

GMRT H_I Observations of the Eridanus Group of Galaxies

A. Omar^{*1} & K. S. Dwarakanath²

Raman Research Institute, Sadashivanagar, Bangalore 560 080, India.

¹*e-mail: aomar@upso.ernet.in*

²*e-mail: dwaraka@rri.res.in*

Received 2004 August 23; accepted 2005 March 22

Abstract. The GMRT H_I 21cm-line observations of galaxies in the Eridanus group are presented. The Eridanus group, at a distance of ~ 23 Mpc, is a loose group of ~ 200 galaxies. The group extends to more than 10 Mpc in projection. The velocity dispersion of the galaxies in the group is ~ 240 km s⁻¹. The galaxies are clustered into different sub-groups. The overall population mix of the group is 30% (E + S0) and 70% (Sp + Irr). The observations of 57 Eridanus galaxies were carried out with the GMRT for ~ 200 h. H_I emission was detected from 31 galaxies. The channel rms of ~ 1 mJy beam⁻¹ was achieved for most of the image-cubes made with 4 h of data. The corresponding H_I column density sensitivity (3σ) is $\sim 1 \times 10^{20}$ cm⁻² for a velocity-width of ~ 13.4 km s⁻¹. The 3σ detection limit of H_I mass is $\sim 1.2 \times 10^7$ M_⊙ for a line-width of 50 km s⁻¹. Total H_I images, H_I velocity fields, global H_I line profiles, H_I mass surface densities, H_I disk parameters and H_I rotation curves are presented. The velocity fields are analysed separately for the approaching and the receding sides of the galaxies. These data will be used to study the H_I and the radio continuum properties, the Tully–Fisher relations, the dark matter halos, and the kinematical and H_I lopsidedness in galaxies.

Key words. Galaxies: groups, clusters—individual: Eridanus—radio lines: H_I 21cm-line.

1. Motivation

Several redshift surveys carried out over the past several decades indicate that galaxies are distributed inhomogeneously in the local Universe. The regions of highest galaxy densities are superclusters and clusters. However, the majority of galaxies in the local Universe are found in less dense regions called groups. According to theories of hierarchical structure formation, clusters are built via mergers of groups. Clusters differ from groups in several aspects. A remarkable difference is observed in the morphological mix and H_I content of the galaxies. Clusters have an enhanced population of both the early type (S0 and E) galaxies and the H_I depleted spirals (Curtis 1918; Hubble &

^{*}Present address: ARIES, Manora peak, Nainital 263 129, Uttarakhand, India.

Humason 1931; Davies & Lewis 1973; Giovanelli & Haynes 1985; Warmels 1988; Cayatte *et al.* 1990; Bravo-Alfaro *et al.* 2000) while groups are populated mainly by H_I rich spirals. Dressler (1980) noticed a tight correlation between the galaxy morphology and the local projected galaxy density. This correlation, known as the density-morphology relation, is observed to be valid over more than five orders of magnitude in the projected galaxy density (Postman & Geller 1984). The origin of the enhanced population of E + S0s in high galaxy density regions has been the subject of much debate. There are two hypotheses for the formation of S0s, one is “Nature” where it is believed that early types were formed as such, and the other is “Nurture” according to which S0s are transformed spirals. Some of the recent observations indicate that the clusters at intermediate redshifts ($z \sim 0.1 - 0.3$) tend to have a higher fraction of S0s at the expense of spirals (e.g., Poggianti *et al.* 1999; Dressler *et al.* 1997; Fasano *et al.* 2000). These observations support the “Nurture” scenario.

Several gas-removal mechanisms, *viz.*, ram-pressure stripping (Gunn & Gott 1972), thermal conduction (Cowie & Songaila 1977), viscous stripping (Nulsen 1982), harassment (Moore *et al.* 1998), starvation, etc. have been proposed to explain the H_I deficiency in cluster spirals. Some of these processes are also believed to be driving the transformation of spirals to S0s. Each of these processes has been predicted to remove H_I mass of the order of the typical H_I mass of a galaxy. These processes, however, are quite sensitive to several parameters like the density of the intra-cluster medium (ICM), the radial velocity vector of the galaxy, the magnetic field in the ICM, the gas-reservoir in the galaxy halo, etc. Some of these parameters have poor estimations which make the efficacy of these processes doubtful. It has been argued that no single gas-removal mechanism can explain the global H_I deficiency in cluster spirals (e.g., Magri *et al.* 1988). The exact physical mechanism(s) responsible for H_I depletion in cluster spirals, therefore, remains uncertain. These difficulties have led one to speculate that cluster galaxies were perhaps H_I deficient even before they fell into the cluster. Such a scenario can be explored by studying groups of galaxies.

Several groups have been previously imaged in H_I, e.g., the Hickson Compact Groups (HCGs; Verdes-Montenegro *et al.* 2001) and the Ursa-Major group (Verheijen & Sancisi 2001). The galaxy densities in HCGs are comparable to that in galaxy clusters, although HCGs have far less number of galaxies compared to that in clusters. The galaxies in some of the HCGs were found to be significantly H_I depleted. HCGs also tend to have a significant population of early type galaxies. The Ursa-Major group, which has only a few S0s and no ellipticals, showed no significant H_I deficiency. The environment in the Ursa-Major group is similar to that in field. Here, we present an H_I survey of the Eridanus group of galaxies with the recently completed Giant Meterwave Radio Telescope (GMRT). The Eridanus group is believed to be at an evolutionary stage intermediate to that of field and a cluster. The Eridanus group has a significant population of early type galaxies. The sub-grouping of galaxies in the group is quite prominent. The Eridanus group also has weak diffuse X-ray emission centered around some of the brightest galaxies in the sub-groups. On a broader perspective, the properties of the Eridanus group are between that of a loose group like the Ursa-Major and a cluster like Fornax or Virgo. The main aim of this survey is to identify the galaxy evolution processes active in an environment intermediate between that of a cluster and field.

The GMRT observations provided both the H_I and the radio continuum ($\nu \sim 1.4$ GHz) data. The kinematical information of galaxies has also been obtained

using the H_I data. This survey has capabilities to carry out several other studies. Some of the studies proposed to be carried out are the following:

- H_I content of galaxies in the Eridanus group.
- H_I morphologies of galaxies in the group.
- Tully–Fisher relations.
- Radio–Far-infrared correlation.
- Rotation curves and dark matter halos.
- Kinematical and H_I lopsidedness.

In the present paper, the GMRT observations and the data analyses are described. We also investigate correlations between H_I and optical properties in this paper. The paper is arranged in the following order. The next section describes the properties of the Eridanus group. Section 3 contains details of the GMRT observations. The analyses of the H_I images are described in section 4. Some of the H_I properties of the Eridanus galaxies are discussed in section 5. The results are presented in the tables in Appendix A. The H_I atlas consists of the H_I images, the H_I velocity fields, the global H_I profiles, the H_I surface densities, the H_I rotation curves, and the kinematical parameters of the H_I disks. The H_I atlases are given in Appendix B.

2. The Eridanus group

2.1 Introduction

The concentration of galaxies in the Eridanus region is known for many decades (Baker 1933, 1936). The complex morphology of this region was pointed out by de Vaucouleurs (1975). The Eridanus group was identified as a moderate size cluster in a large scale filamentary structure near $cz \sim 1500 \text{ km s}^{-1}$ in the Southern Sky Redshift Survey (SSRS; da Costa *et al.* 1988). This filamentary structure, which is the most prominent in the southern sky, extends for more than 20 Mpc in projection. The Fornax cluster and the Dorado group of galaxies are also part of this filamentary structure. The dynamical parameters of a few galaxies in the Eridanus group were first published by Rood *et al.* (1970). With the increased number of identifications in this region by Sandage & Tammann (1975) and Welch *et al.* (1975), the latter authors speculated a dynamical connection between the Fornax cluster and the Eridanus group. Using the data from the Southern Galactic Cap sample (SGC; Pellegrini *et al.* 1990), Willmer *et al.* (1989) grouped the galaxies in the Eridanus region into different sub-groups and studied their dynamics. They concluded that each sub-group is a bound structure and possibly the entire group is also gravitationally bound with a dynamical mass greater than $10^{13} M_{\odot}$. They further pointed out that the Fornax and the Eridanus together constitute a bound system. The Eridanus group is a dynamically young system with a velocity dispersion of $\sim 240 \text{ km s}^{-1}$, which is lower compared to that ($\sim 1000 \text{ km s}^{-1}$) seen in clusters like the Coma. The distance to the group is estimated as $23 \pm 2 \text{ Mpc}$ based on the surface brightness fluctuation measurements (Tonry *et al.* 1997; Jensen *et al.* 1998; Tonry *et al.* 2001). All identified members in the group are in the Heliocentric velocity range of $\sim 1000\text{--}2200 \text{ km s}^{-1}$, except NGC 1400 (S0), which has a velocity of $\sim 558 \text{ km s}^{-1}$. However, NGC 1400 is predicted to be at a similar distance as that of the other members of the group based on the surface brightness fluctuation measurements.

2.2 Group structure, membership, and morphological mix

The velocity-cone diagrams are plotted in Fig. 1. The plots are in the supergalactic coordinates. The velocities obtained from the NASA Extra-galactic Database (NED) are Heliocentric, and follow the optical definition. The clustering of galaxies near $l = 283^\circ$ and $b = -43^\circ$ is the Eridanus group. Most of the galaxies are concentrated in the velocity range $cz = 1000\text{--}2200 \text{ km s}^{-1}$. The group appears to be loose and irregular. The clustering of galaxies near the apex ($cz = 0$) is the local group. In Fig. 2, the positions of galaxies within the velocity range $500\text{--}2500 \text{ km s}^{-1}$ are plotted. There are 181 galaxies in this plot, 60 early types (E + S0) and 121 late types (Sp + Irr). The approximate boundaries of three main sub-groups identified by Willmer *et al.* (1989) are marked in this figure. The sub-clustering of galaxies can be seen in this figure.

It can be seen that most of the early type galaxies are in the region inside the circle marked in Fig. 2. The sub-clustering is also prominent in the inner region. In the outer regions, population is dominated by spirals. The morphological mix is appreciably different in each sub-group. The three sub-groups namely NGC 1407, NGC 1332, and NGC 1395 have their brightest members as an elliptical or an S0. The NGC 1407 sub-group is the richest in the early types, most of them being S0s. The population of (E + S0) and (Sp + Irr) in the NGC 1407 sub-group is 70% and 30% respectively, while that in most of the other sub-groups is $\sim 40\%$ and 60% respectively. The overall population mix of the Eridanus group is $\sim 30\%$ (E + S0) and 70% (Sp + Irr). The velocity-histograms of the early types and the late types are plotted in Fig. 3. There is no appreciable difference in the velocity range over which the early types and the late types are distributed. However, it can be seen from the upper panel of Fig. 3 that the distribution of early-type galaxies is approximately a Gaussian whereas the late-type galaxies are rather uniformly distributed. Upon inspection of the locations of late-type galaxies at lower ($1000\text{--}1300 \text{ km s}^{-1}$) and higher ($1800\text{--}2100 \text{ km s}^{-1}$) velocity ends from the mean, it appears that the galaxies having higher velocities are more uniformly distributed in the sky compared to those at lower velocities. The lower velocity galaxies are largely confined within the circle drawn in Fig. 2. Further, it is interesting to note that the population mix of the NGC 1407 sub-group is similar to that seen in evolved clusters like the Coma, whereas the velocity dispersion ($\sim 250 \text{ km s}^{-1}$) of the NGC 1407 sub-group is much smaller than that of Coma ($\sim 1000 \text{ km s}^{-1}$). Some other groups like Leo I, NGC 3607, and NGC 5846, which have lower velocity dispersions compared to that in clusters are also populated mainly by early type galaxies.

2.3 X-ray properties

The optically bright early type galaxies NGC 1400, NGC 1407, NGC 1395, and NGC 1332 are known X-ray sources in this group. Trinchieri *et al.* (2000) reported the presence of diffuse X-ray emission around NGC 1407. The processed and calibrated X-ray images ($0.1 \text{ keV--}2.0 \text{ keV}$) centered at NGC 1407, NGC 1395, and NGC 1332 were obtained from the ROSAT PSPC (Roentgen Satellite Position Sensitive Proportional Counter) archival data. Each field was observed for $\sim 6 \text{ h}$ using the ROSAT PSPC instruments. The soft X-ray images shown in Fig. 4 were convolved with a circular Gaussian beam of FWHM $90''$ to enhance the diffuse emission. Apart from the

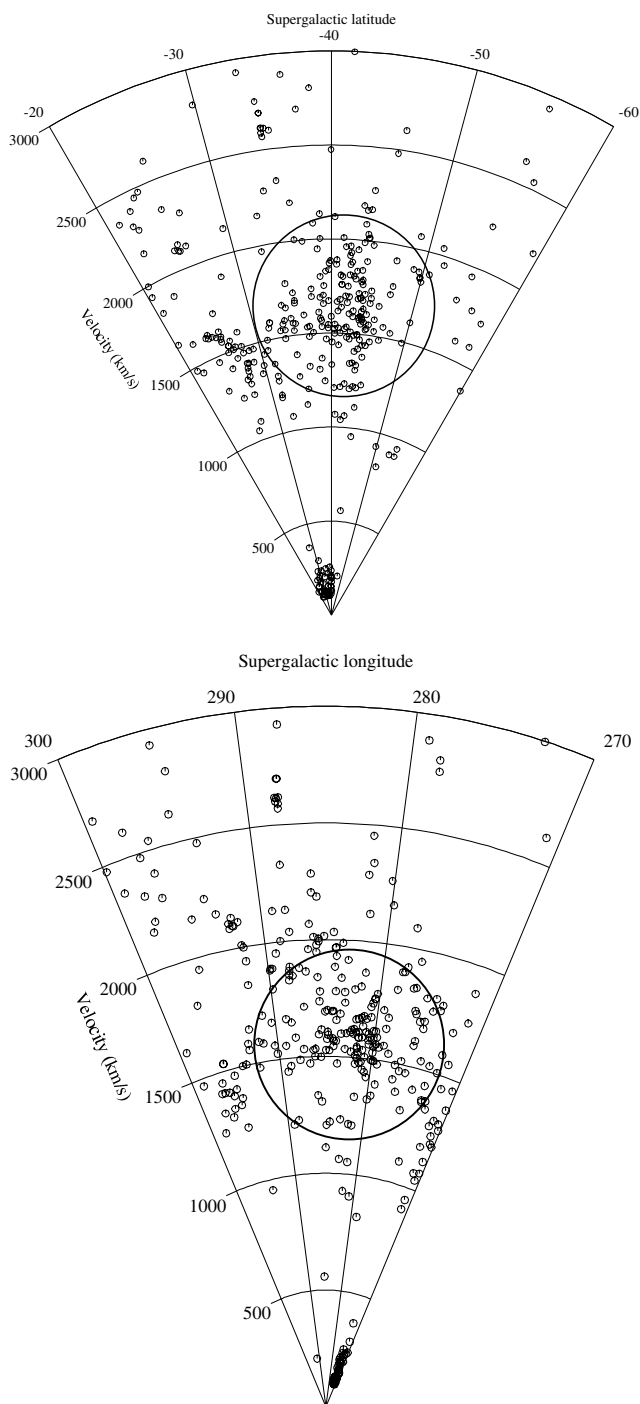


Figure 1. Positions of galaxies in the velocity-cone diagrams. The velocities are Heliocentric from the NASA Extra-galactic Database (NED). The circles mark the Eridanus group.

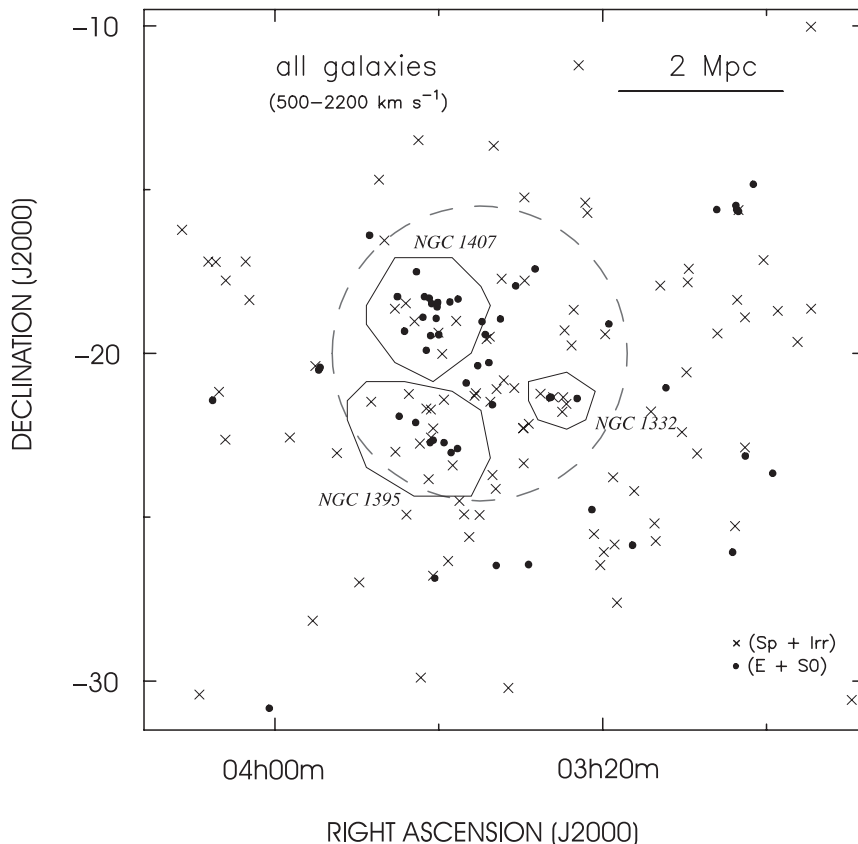


Figure 2. Galaxies in the Eridanus group. The radius of the dashed circle is ~ 2 Mpc within which galaxies were observed with the GMRT. A few sub-groups are marked with their approximate boundaries as identified by Willmer *et al.* (1989).

X-ray emission associated with NGC 1400, NGC 1407 and a few other unresolved sources in the field, diffuse emission centered at NGC 1407 and NGC 1395 can be seen in Fig. 4. The extent of the diffuse emission is $\sim 30'$ (~ 200 kpc) around NGC 1407 and $\sim 20'$ (~ 135 kpc) around NGC 1395. No diffuse emission is seen around NGC 1332 (not shown here). Using the PIMMS (Portable, Interactive Multi-Mission Simulator; Mukai 1993) tool, the diffuse emission was modeled as thermal free-free emission from a Raymond-Smith plasma of energy ~ 1.0 keV ($T \sim 10^7$ K) and metallicity 0.2-solar. The choice of the temperature and the metallicity is in accordance with typical values found in X-ray groups (Mulchaey 2000).

The total X-ray luminosity of the diffuse emission in the energy range 0.1–2.0 keV is 1.6×10^{41} erg s $^{-1}$ for the NGC 1407 sub-group and $\sim 6.8 \times 10^{40}$ erg s $^{-1}$ for the NGC 1395 sub-group. The intra-group medium density is estimated as $\sim 2.0 \times 10^{-4}$ cm $^{-3}$ in the X-ray emitting region. The X-ray luminosity of the Eridanus group is about 2–3 orders of magnitude lower compared to that of the clusters like Coma and Virgo. The estimated intra-group medium density is about an order of magnitude lower than that observed in virialised clusters like Coma.

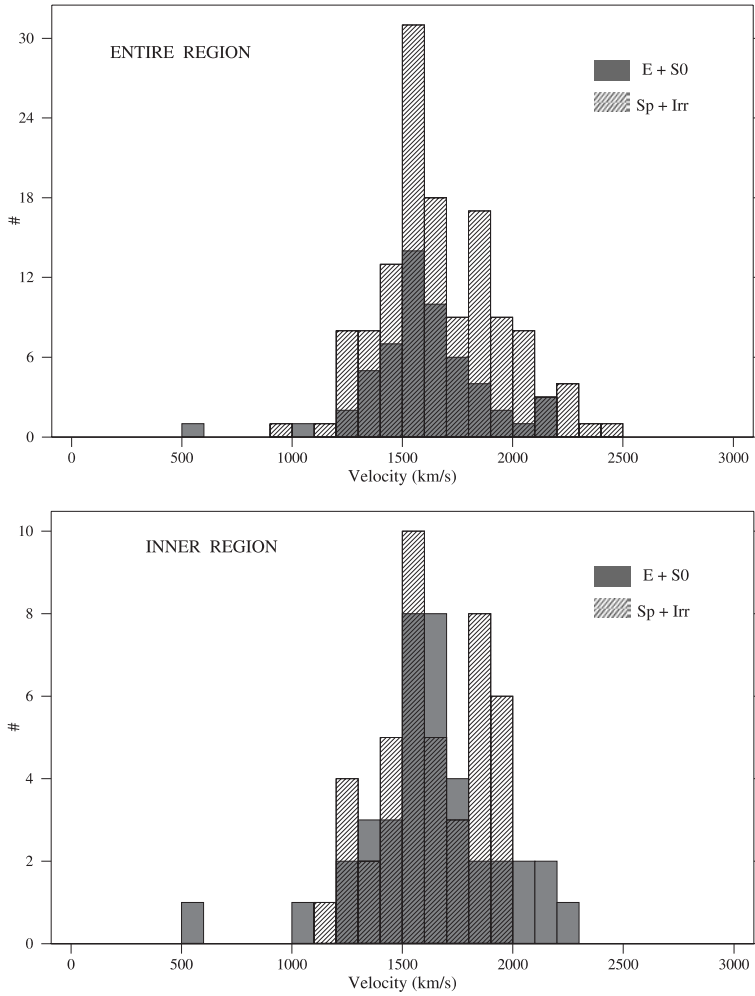


Figure 3. Velocity-histograms of galaxies in the inner 4 Mpc region (bottom panel) and in the entire group (top panel) are shown. There is no appreciable difference in the velocity range over which the early types and the late types are distributed.

2.4 Comparison of the Eridanus group with other groups and clusters

The properties of the Eridanus group are compared with the Virgo and the Fornax cluster, and the Ursa-Major group in Table 1. All these systems are at comparable distances. Both the Fornax cluster and the Eridanus group belong to the filamentary structure described by da Costa *et al.* (1988). The Ursa-Major group is a loose group of galaxies (Tully *et al.* 1996). All the four systems have quite different properties. The Fornax cluster having the highest galaxy density has the lowest spiral fraction, consistent with the density–morphology relation. The Eridanus group is intermediate between the Virgo cluster and the Ursa-Major group in terms of its velocity dispersion, its X-ray luminosity, and its number of early-type galaxies. The mean projected galaxy density in Eridanus is intermediate between that in Ursa-Major and in Virgo. The

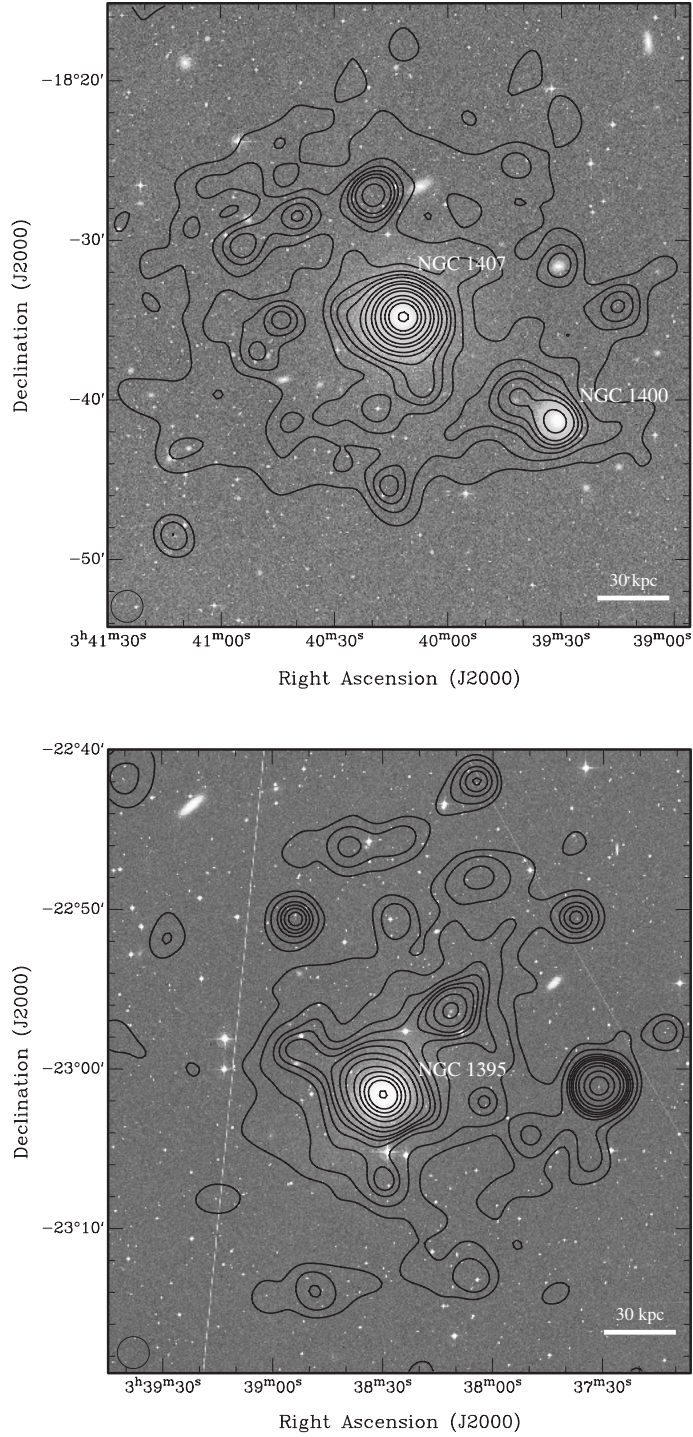


Figure 4. Contours of X-ray emission around NGC 1407 and NGC 1395 overlaid upon the optical images from the DSS. The X-ray images are retrieved from the ROSAT PSPC archived data and smoothed with a circular Gaussian beam of $90''$.

Table 1. Comparison of four nearby galaxy groups and clusters.

Properties	Virgo ^a	Fornax ^b	Eridanus ^c	Ursa-Major ^d
Distance (Mpc)	17	20	23	21
No. of E + S0s	71	23	36	9
No. of Sp + Irr's	123	17	42	53
(S + Irr) fraction	0.6	0.4	0.5	0.9
Vel. dispersion (km s ⁻¹)	760	350	240	150
log X-ray luminosity (erg s ⁻¹)	43.5	41.7	41.4	–
References	1,2,3,4	2,5,6,7	8	9,10

Notes – **(a)**: Inner 6° region, **(b)**: Inner 2°.4 region, **(c)**: Inner 9° region, **(d)**: Inner 15° region. **References** – (1) Federspiel *et al.* (1998), (2) Ferguson (1989), (3) Binggeli *et al.* (1987), (4) Mushotzky & Smith (1980), (5) Mould *et al.* (2000), (6) Richter & Sadler (1985), (7) Paolillo *et al.* (2001), (8) Omar (2004), (9) Sakai *et al.* (2000), (10) Tully *et al.* (1996).

galaxies in Virgo are H_I deficient. The Ursa-Major group has normal H_I content. The X-ray luminosity of the Eridanus group is at the lower end of the X-ray luminosities observed in groups (Mulchaey 2000). The velocity dispersion of the galaxies in the Eridanus group is intermediate between that in Fornax and in Ursa-Major. From this comparison, it appears that the Eridanus group forms a system which is intermediate between a loose group (Ursa-Major) and a rich cluster (Virgo and Fornax).

3. Observations and data reduction

The present GMRT H_I observations offer several advantages over studies carried out in the past using single dish telescopes. The GMRT is an interferometric array of thirty 45-m diameter fully steerable parabolic dishes. A description of the GMRT is given by Swarup *et al.* (1991). The GMRT is located at a site (longitude = 74°.05 E, latitude = 19°.092 N, altitude ~ 650 m above MSL) about 80 km north of Pune, India. The configuration of the GMRT is optimized to meet the requirements of high angular resolution and of being able to image extended emission. This optimization is achieved through a hybrid configuration of the antennas. Fourteen of the thirty dishes are located more or less randomly in a compact central array within an area of about 1 × 1 square kilometer, and the remaining sixteen dishes are spread out along the 3 arms of an approximately Y-shaped configuration over a larger region. The longest separation of antennas is ~ 25 km, and the shortest separation is ~ 100 m. The GMRT is expected to be sensitive to structures on the scales of 2''–7' at a wavelength of 21 cm. The angular sizes of the galaxies in the Eridanus group are in the range 1'–5' implying that the data should be sensitive to image radio emission (H_I and continuum) over the full extents of galaxies. The FWHM of the primary beam of a GMRT antenna is ~ 24' at 1.4 GHz.

3.1 Sample of galaxies

The selection of galaxies for the H_I observations were made keeping in mind the broad perspective of the work. The galaxies were not selected based on their H_I contents or

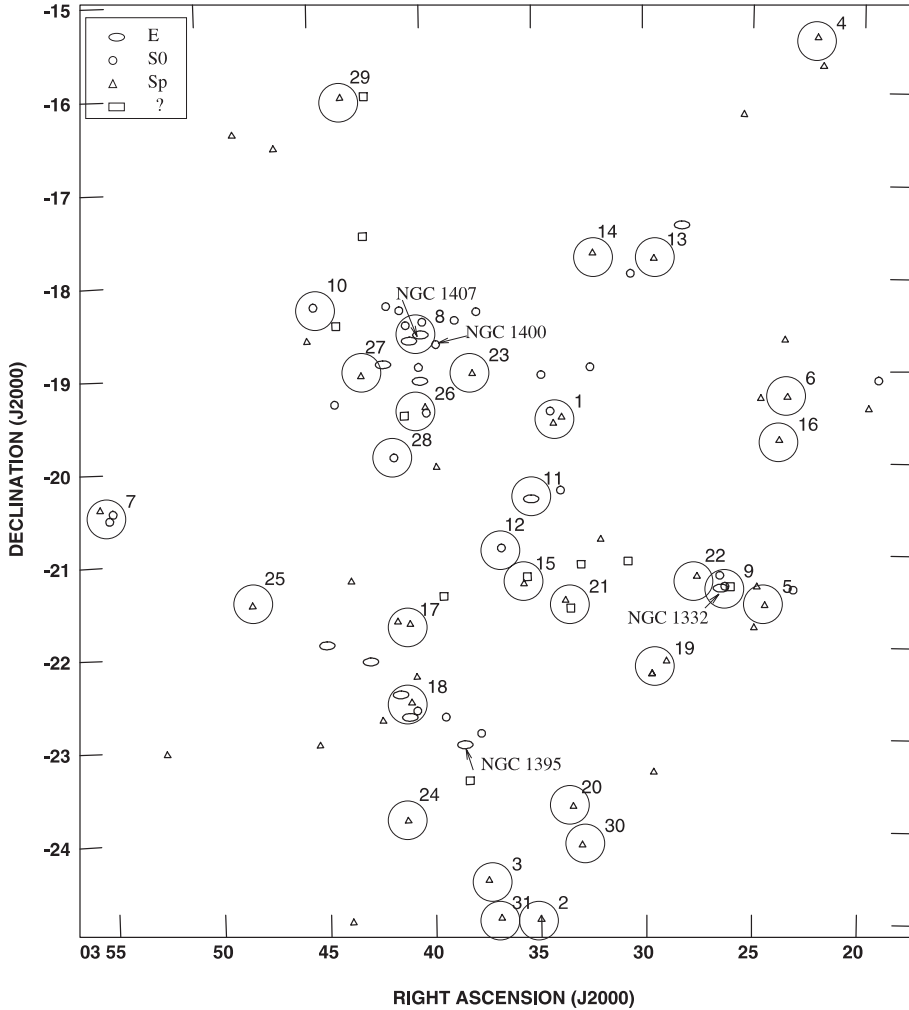


Figure 5. The fields observed with the GMRT. The bigger circles correspond to the FWHM of the GMRT primary beam ($\sim 24'$) at 1.4 GHz.

their optical luminosities. Both early type and late type galaxies were included in the sample. The galaxies were selected from the inner 4 Mpc region of the group where galaxy density is higher and most of the S0s are found. A follow up R-band photometric observations were also carried out with the 1-m optical reflector at the Aryabhata Research Institute of Observational Sciences (ARIES; *formerly* State Observatory), Nainital. The optical data analysis is presented in Omar (2004).

Since the present H_I study with the GMRT was carried out with a limited telescope time of ~ 200 h, the pointing centres of the observations were adjusted in a way to include two or more galaxies within the FWHM of the primary beam. Unfortunately, one complete run of observations on 16 galaxies (during November 2001), mostly early types, was badly affected due to ionospheric scintillations, perhaps related to the intense solar activities during that year. The data collected during this period could not be used to obtain images. Five galaxies from these lost observations were re-observed

later in 2002. The science quality data were obtained for a total of 46 galaxies. In Tables 2 and 3, the complete observed sample of 57 galaxies is listed with some of their previously known optical and radio properties.

3.2 Observational parameters

The Eridanus group can be observed with the GMRT for ~ 8 h in a given day. Often, two galaxies were observed in each day. The observing strategy was optimized to get uniform distribution of visibilities. Two galaxies were observed alternately for 15–20 min each followed by 5–7 min of observations of secondary calibrators. This cycle was repeated and a total of 3–4 h of observing time was accumulated on each galaxy. Most of the observations were carried out using an 8 MHz bandwidth over 128 channels, which gives a velocity resolution of $\sim 13.4 \text{ km s}^{-1}$. The observations were carried out for longer duration (~ 8 h) for some of the early type galaxies, and with smaller bandwidths (2–4 MHz) for smaller inclination galaxies to get sufficient velocity resolution. A total of ~ 200 h of the GMRT observations were carried out spread over 3 years (2000–2002). The data obtained during November 2001 which were corrupted due to scintillations were discarded. The observing parameters are listed in Table 4.

The VLA calibrators 0240-231 and 0409-179 were used as the secondary calibrators. 0240-231 is classified as “un-resolved” for all the four VLA configurations with a 20-cm flux density of 6.3 Jy. 0409-179 is resolved by baselines longer than 10 km with a 20-cm flux density of 2.2 Jy. 0137 + 331 (3C 48) and 0542 + 498 (3C 147) were used as the primary calibrators. 3C 48 is resolved by baselines longer than 8 km with a 20-cm flux density of 16.5 Jy and 3C 147 is resolved by baselines longer than 10 km with a 20-cm flux density of 22.5 Jy. 3C 48 was observed in the beginning and 3C 147 was observed at the end of each observing run for 20–30 min. The flux densities of the primary calibrators were estimated at the observed frequencies using their known radio spectra from the VLA observations in the 1999.2 epoch.

3.3 Data acquisition and reduction

The data (visibilities) were collected in the LTA (Long Time Accumulation) format, which is the native format for the GMRT data. The LTA data were converted to FITS (Flexible Image Transport System) format for subsequent processing. Visibilities were averaged over ~ 16 s. The data were monitored online. The data were later flagged from the antennas having low gains, for time ranges where data were corrupt, and at lower elevations (usually below 25°) where correlation drops significantly (below 50% in some cases). The flux densities of the secondary calibrators were estimated based on the flux densities of the primary calibrators. The visibilities on the target galaxies were calibrated by interpolating the complex gains determined using the secondary calibrators. Since the spectral responses of filters are not flat, the initial calibration was carried out using the data averaged over four to six channels. The spectral response of the antennas were determined using both the secondary and the primary calibrators, and an averaged spectrum was used to correct the band shapes. The gains start declining significantly after the 110th channel. The first 1–3 channels are generally corrupted in the filter response. Therefore, the data were used between channels 3 and 115. An initial $H\text{I}$ spectrum was generated using the AIPS (Astronomical Image Processing System) task POSSM at spatial frequencies below $2 \text{ k}\lambda$ in the direction

of target galaxies. This range of spatial frequencies enables most of the H_I signals to be detected in the H_I spectrum. This spectrum is to identify channels with H_I emission.

The continuum-data were generated by averaging the channels devoid of H_I line emission. The continuum images were made using this channel averaged data and were used for self-calibration. Several iterations of phase self-calibrations were performed to improve the dynamic range of the images. The final self-calibrated solutions were applied to the line-data. The continuum emission was subtracted from the line-data using the AIPS tasks UVSUB and UVLIN. The self-calibrated and continuum subtracted line-data were used to make the image cubes at different resolutions by selecting appropriate (u, v) ranges. The image cubes were made at two resolutions – one with a resolution of $\sim 15''$ (high resolution cube) using (u, v) data in the range $0.2\text{--}20\text{ k}\lambda$, and another with a resolution of $\sim 50''$ (low resolution cube) using (u, v) data in the range $0.2\text{--}5\text{ k}\lambda$.

4. Image analysis

The images were analysed using the GIPSY (Groningen Image Processing System) package developed by the Kapteyn Institute, the KARMA visualization tool (Gooch 1996), and the AIPS package developed by the National Radio Astronomy Observatory.

Since the angular resolution varied by a few arc second in different cubes, all high resolution cubes were convolved to a common resolution of $20'' \times 20''$. In some cases, intermediate resolution cubes at $25''$ or $30''$ were also made. The channel images typically have an rms of 1 mJy beam^{-1} . The 3σ column density detection limit in the $20''$ images is $1 \times 10^{20}\text{ cm}^{-2}$. The cubes are sensitive (3σ) to detect a galaxy of H_I mass $1.2 \times 10^7 M_{\odot}$ for an H_I line-width of 50 km s^{-1} . The image cubes were inspected visually to identify H_I signals. The channel images are presented for all the H_I detected galaxies elsewhere (Omar 2004). An example of the channel images is shown in Fig. 6.

4.1 Total H_I map and H_I diameter

The zeroth and the first order moment maps were generated at both the low ($50''$) and the high ($20''$) angular resolutions. The moment zero map or the total H_I image is obtained by summing the H_I images in different channels. The cubes were first blanked to separate the H_I signals from noise before summing the channels. The blanking can be done in several ways. The total H_I image depends on the blanking procedure (Rupen 1999). One of the methods is to blank the pixels below a certain flux density level. A higher cutoff (e.g., 5σ) makes the total H_I images patchy while a lower cutoff (e.g., 3σ) makes the images noisy. The low surface brightness nature of H_I emission makes it difficult to separate the low level signals from noise. A hybrid approach has been shown to be effective in overcoming this problem (Rupen 1999). This approach involves masking the noise after smoothing the cube using the AIPS task MOMNT. The moment maps are still estimated using the un-smoothed cube in this approach.

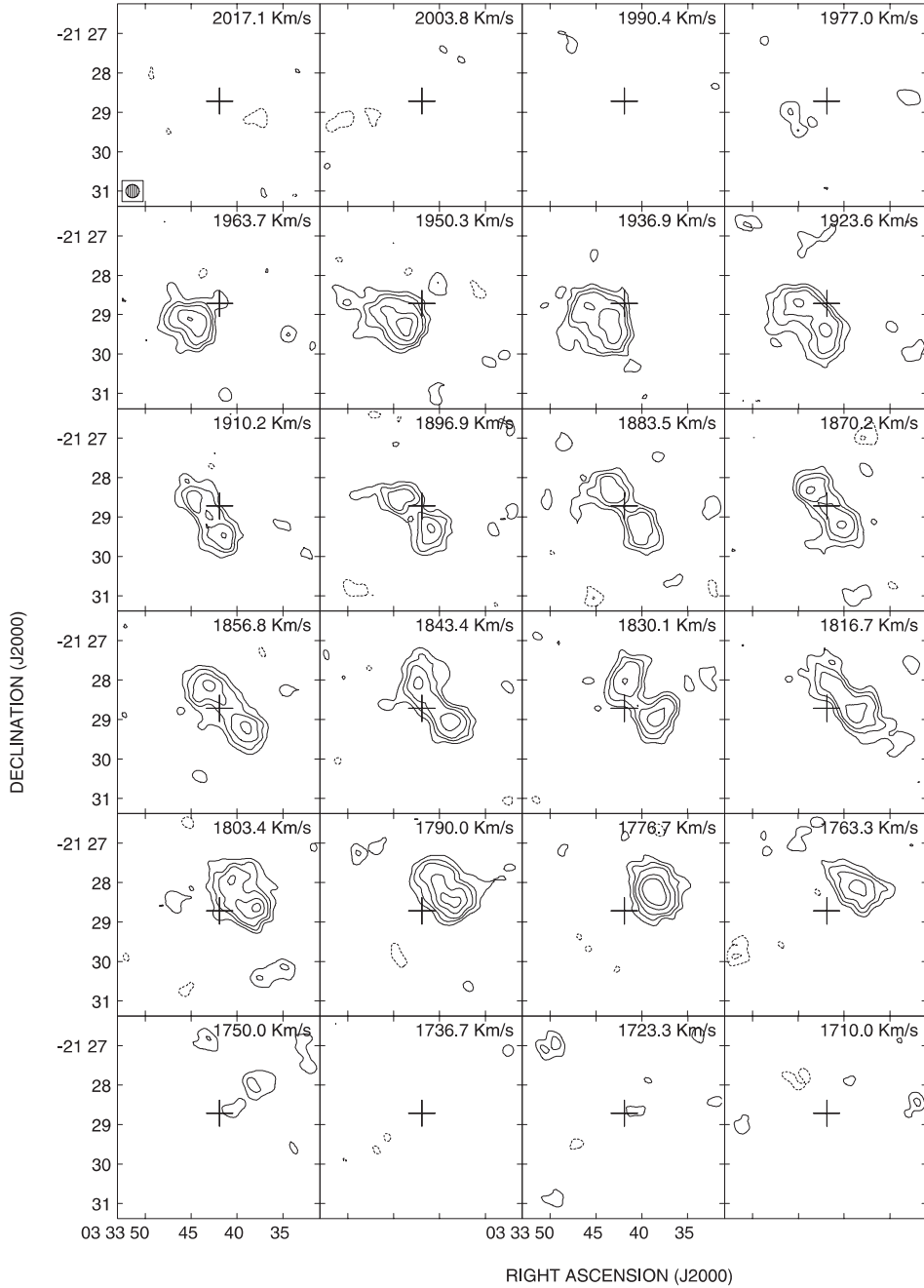


Figure 6. $H\text{I}$ emission from IC 1953. The ‘+’ sign marks the optical centre of the galaxy. The rms/channel is $1.4 \text{ mJy beam}^{-1}$. The contours are at 2.5, 3.75, 5, 7.5, and 10 times the rms. The images are convolved with a circular Gaussian beam of $20'' \times 20''$.

The flux density (mJy beam^{-1}) is converted to H I column density using the following relation (equation 3.38, Spitzer 1978):

$$N_{\text{H I}}(\alpha, \delta) = \frac{1.1 \times 10^{21} \text{ cm}^{-2}}{\theta_a \times \theta_b} \delta v \sum_{j=1}^{N_{\text{chan}}} S_j(\alpha, \delta), \quad (1)$$

where θ_a and θ_b are the FWHM of the synthesised beam measured in arc second along the major and minor axes respectively. S_j is the H I flux density (mJy beam^{-1}) in the channel j and δv is the velocity resolution in km s^{-1} . The H I gas is assumed to be optically thin. A collage of the integrated H I maps of the Eridanus galaxies is shown in Fig. 7 as contours and in Fig. 8 as colour-coded. The individual galaxies are enlarged ten times. To avoid overlap, some galaxies are slightly displaced from their actual positions.

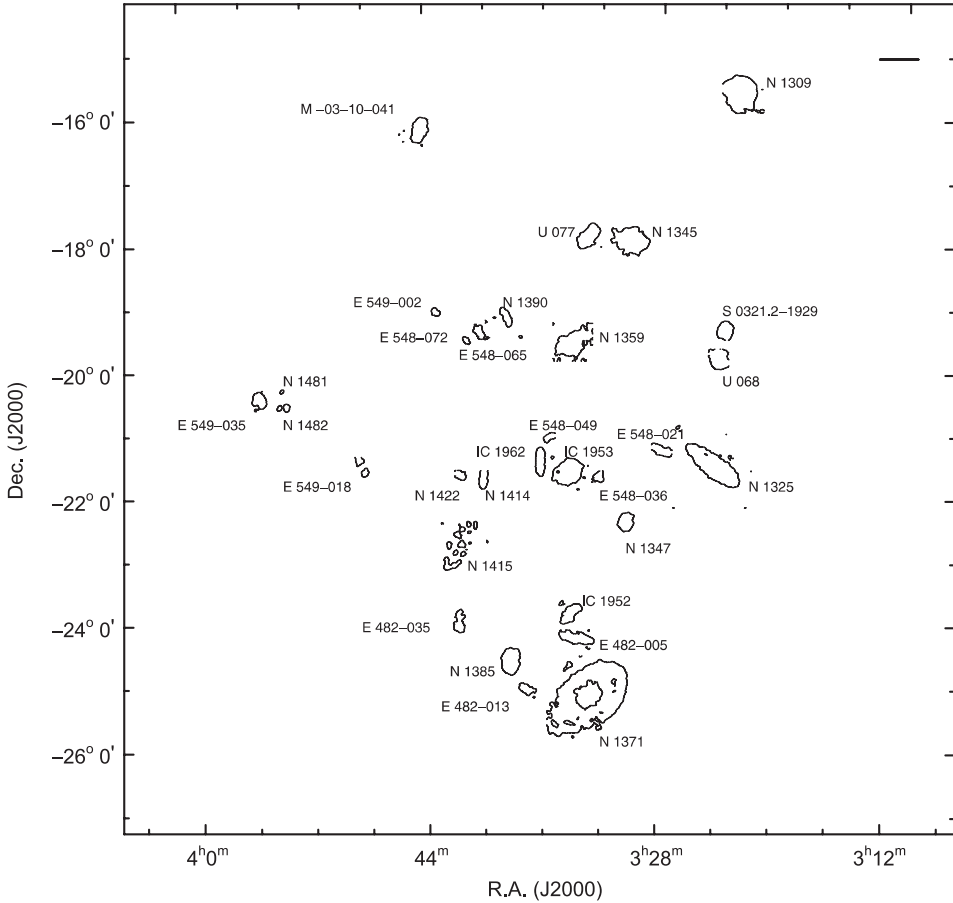


Figure 7. A contour-image collage of the H I detected galaxies in the Eridanus group. Only one contour is plotted to indicate the extent of each galaxy at $N_{\text{H I}} = 10^{20} \text{ cm}^{-2}$. The individual galaxies are magnified ten times. To avoid overlap, some galaxies are slightly displaced from their actual positions. A bar at the upper right hand corner indicates a scale of 20 kpc for the enlarged sizes of the galaxies. Otherwise, 1° corresponds to $\sim 400 \text{ kpc}$.

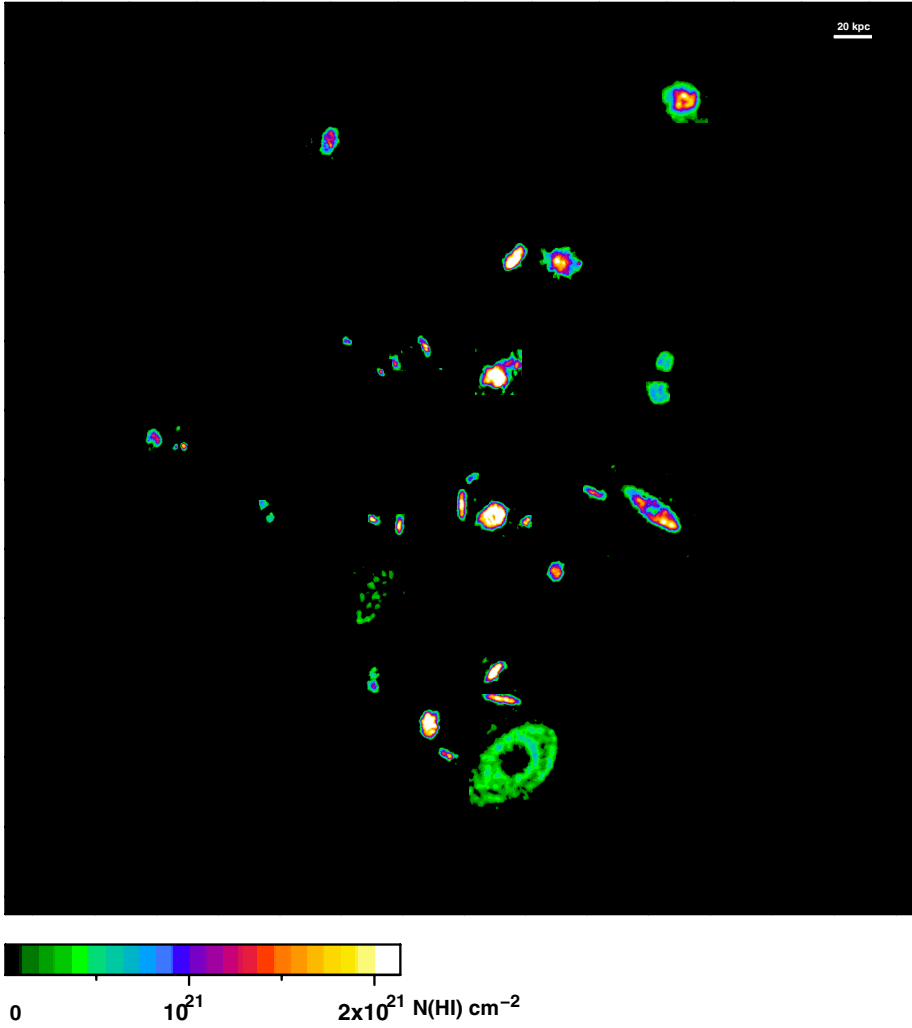


Figure 8. A collage of the $H\text{I}$ detected galaxies in the Eridanus group. The colour key indicates the $H\text{I}$ column density.

The diameters of the $H\text{I}$ disks were estimated from the high resolution total $H\text{I}$ images at a fixed face-on $H\text{I}$ surface density of $1 \text{ M}_{\odot} \text{ pc}^{-2}$. Due to projection effects, the sensitivities to the face-on $H\text{I}$ surface densities were not uniform. Therefore, in some cases the $H\text{I}$ diameters were extrapolated to the face-on $H\text{I}$ surface density level of $1 \text{ M}_{\odot} \text{ pc}^{-2}$.

4.2 $H\text{I}$ velocity field

The conventional way of deriving the velocity field is to compute the intensity-weighted first order moment of the $H\text{I}$ images at different velocities. There is an alternative to obtain velocity fields by fitting the $H\text{I}$ profiles at every pixel with a Gaussian. These

profiles are usually asymmetric depending upon the kinematics of H_I along the line-of-sight, and also due to the beam smearing caused by the finite size of the synthesised beam. The broadening and the asymmetry will depend upon the H_I distribution in the galaxy. The effect of beam smearing will be more pronounced in edge-on systems. As a result of the asymmetry and the broadening in the H_I profile, a single Gaussian component will not give an accurate result. Unfortunately, multi-component Gaussian fit could not be carried out as the signal-to-noise ratio of the detections were not sufficient.

Rupen (1999) has briefly discussed the merits and drawbacks of these two procedures. In the present analysis, the first order moment maps were found to be generally noisier than the velocity field maps obtained by Gaussian fits. This may be because the Gaussian fits were not as sensitive to the outliers as moment maps were. Therefore, in the present analysis, Gaussian fits were used to construct the velocity field maps. It should be noted that both the procedures to obtain the velocity field will underestimate rotation velocities at locations of steep velocity gradients. The flat part of the rotation curve, however, remains unaffected.

4.3 Global H_I profiles

The integrated H_I flux density as a function of velocity is the global H_I profile as would have been obtained from a single dish observation. The low resolution (50'') cubes were used to obtain the global profiles as these cubes are most sensitive to the diffuse emission. The lower resolution zeroth order moment map was used to mark the region over which the flux density was estimated in the channel images. The H_I mass is obtained by using the following relation:

$$M_{\text{H I}} (M_{\odot}) = 2.36 \times 10^5 D^2 \delta v \sum_{j=1}^{N_{\text{chan}}} S_j \quad (2)$$

where D is the distance in Mpc, S_j is the integrated flux in Jy in the spectral channel j of velocity width δv in km s^{-1} . The distance is taken as 23 Mpc. In Fig. 9, the integrated H_I flux ($\delta v \sum_{j=1}^{N_{\text{chan}}} S_j$) from the GMRT is compared with that from the HIPASS data and from other single dish data. Some of the galaxies with higher values of the integrated H_I flux in the single dish data show significantly less flux in the GMRT. The H_I disk sizes of these galaxies are among the largest ($> 6'$) in our sample. We believe that the loss of flux for the large galaxies in the GMRT images is due to inadequate sampling of shorter (u, v) spacings in the GMRT.

4.4 H_I line-width

The H_I global profiles often peak at the two extreme ends of the rotation velocities of galaxies. The detailed shape of an H_I profile depends on the rotation curve, the inclination and the H_I distribution in the galaxy. The H_I line-widths are broadened from their true values due to random motions in the H_I gas and due to the finite spectral resolution. The H_I line-width is a crucial parameter for studying the Tully–Fisher (TF) relation. Most of the TF studies still use the H_I line-widths obtained from single dish observations because of the simplicity of such observations. The synthesis data of the Eridanus group of galaxies provide an opportunity to compare the corrected H_I

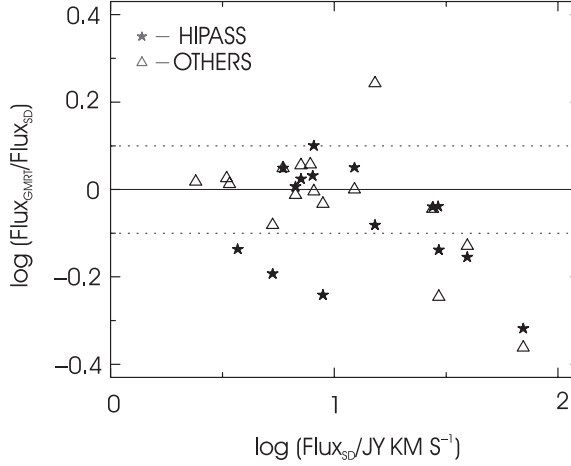


Figure 9. A comparison of the integrated H I flux densities of Eridanus galaxies from the GMRT with those from the single dish data published elsewhere and from the HIPASS. Most of the ratios are within $\pm 25\%$ of unity (indicated by the dotted lines).

line-widths obtained from the single dish H I profiles with those obtained from the H I rotation curves.

For those cases in which double-peaked H I profiles are seen, H I line-widths were estimated at 20% (W_{20}) and at 50% (W_{50}) levels of the peak intensities at the two ends of the H I profiles. The locations of the two peak intensities were estimated separately using Gaussian fits to the profiles. Bottinelli *et al.* (1990) derived an empirical relation to correct for the instrumental broadening. They convolved the H I profile progressively with coarser velocity resolutions for a model galaxy, and determined the broadening. A linear relationship between the channel resolution and the instrumental broadening was suggested. The broadening correction is estimated as $\delta W = 0.55 \times \delta V_i$ for W_{20} and $\delta W = 0.13 \times \delta V_i$ for W_{50} for an instrumental resolution of δV_i . For the current observations, $\delta V_i = 13 \text{ km s}^{-1}$, implying that the corrections are $\sim 7 \text{ km s}^{-1}$ for W_{20} and $\sim 2 \text{ km s}^{-1}$ for W_{50} .

A linear summation of the rotation velocity and the random velocity is appropriate to estimate the observed width for the cases where the intrinsic width is almost boxy (i.e., in fast rotating galaxies). However, a summation in quadrature will be required for the slow rotating (e.g., dwarf) galaxies where the solid body rotation together with the radial distribution of the H I gas will lead to an almost Gaussian profile. A composite relation for all galaxies was given by Tully & Fouque (1985). According to their relation, the width due to the rotation motion W_R , the width due to random and turbulent motions W_t , and the observed width W_l are related by:

$$W_l^2 = W_{R,l}^2 - W_{t,l}^2 \left(1 - 2e^{-(W_l/W_{c,l})^2} \right) + 2W_l W_{t,l} \left(1 - e^{-(W_l/W_{c,l})^2} \right), \quad (3)$$

where the subscript l refers to the level (20%, or 50%) at which the widths are estimated. The $W_{t,l}$ is estimated as $2k_l\sigma$ from the velocity dispersion of the H I gas (σ) due to random and turbulent motions. The constant factor k_l is 1.80 at 20% level and

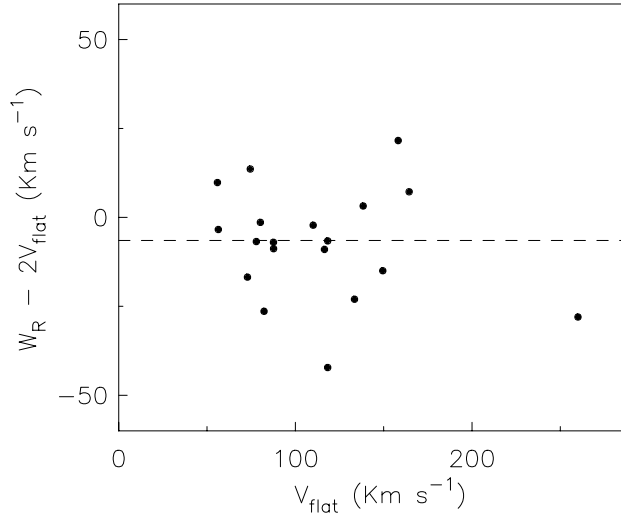


Figure 10. Comparison of the corrected H I widths ($W_{R;50}$) with the flat rotation velocities of the Eridanus galaxies. The dashed line is the mean value at 6.5 km s^{-1} .

1.18 for the 50% level for a Gaussian profile. The value of σ is taken as 6 km s^{-1} . $W_{c,l}$ is a parameter which defines the transition region from linear to quadratic sum. The equation (3) does a linear subtraction if $W_l > W_{c,l}$ and a quadratic subtraction if $W_l < W_{c,l}$. The values of $W_{c,l}$ are determined empirically as 120 km s^{-1} for the 20% level and 100 km s^{-1} for the 50% level by Tully & Fouque (1985).

The corrected H I widths are compared with the flat rotation velocities of the galaxies in Fig. 10. The mean value of the difference ($W_{R,50} - 2V_{\text{flat}}$) is $\sim 6.5 \text{ km s}^{-1}$.

4.5 Rotation curves

The rotation curves were derived using the tilted ring model (cf. Begeman 1989). The GIPSY task ROTCUR was used. The basic methodology of this model is the following. The model assumes the gas to be in circular orbits. The position angle and the inclination of the H I disk are allowed to vary with radius. The fitting procedure generally involves estimation of 5 unknowns, *viz.*, the dynamical centre (X, Y), the systemic velocity (V_{sys}), the position angle (PA) of the major axis, the inclination angle ($INCL$), the circular rotation velocity (V_{rot}), and optionally the expansion velocity V_{exp} . The observed radial velocity $V(x, y)$ at a rectangular sky coordinate (x, y) and at a radius r is given by:

$$V(x, y) = V_{\text{sys}} + V_{\text{rot}} \cos(\theta) \sin(INCL) + V_{\text{exp}} \sin(\theta) \sin(INCL), \quad (4)$$

where

$$\cos(\theta) = \frac{-(x - X) \sin(PA) + (y - Y) \cos(PA)}{r}, \quad (5)$$

$$\sin(\theta) = \frac{-(x - X) \cos(PA) + (y - Y) \sin(PA)}{r \cos(INCL)}. \quad (6)$$

The V_{exp} term was not fitted in the present analysis, and was kept fixed at zero. V_{exp} can be used to estimate the non-axisymmetry in the velocity field. The velocity fields at $20''$ resolution obtained via Gaussian fits were used in this analysis. The iterative procedure described by Begeman (1989) is used to estimate the disk's kinematical parameters and the rotation curve. The velocities are averaged in elliptical annuli of width $10''$ and extracted as a function of the azimuth at different radii in increments of $10''$. The velocity fields in the receding side and in the approaching side were fitted separately to obtain two rotation curves. The iterative scheme to estimate the different parameters is described below.

4.5.1 Dynamical centre (X, Y) and systemic velocity (V_{sys})

The velocity determined from the H I width or the optical velocity of the galaxy was used as the initial guess for the systemic velocity. The guesses for the centre, the position angle and the inclination angle were their respective estimates obtained from the ellipse fits to the optical isophotes. The first iteration is started by fixing the inclination and the position angle, and fitting the centre and the systemic velocity. If the velocity field is symmetric, the fitted values of the dynamical centre and the systemic velocity should be similar for all the rings. However, often galaxies do not have symmetric velocity fields due to kinematical lopsidedness. Therefore, slightly different estimates of centre and systemic velocity may result from each ring. An overlay of the velocity field contours over the optical image helps in deciding the quality of the fit. For a galaxy with no warp, the velocity field lines should run straight along the minor axis. The line joining the cusps of the iso-velocity contours of identical rotation velocity at the two halves of the galaxy should trace the direction of the major axis. The intercept of the major axis with the minor axis is expected to be the dynamical centre.

If a galaxy has a warp, the velocity field lines will show a characteristic integral sign shaped structure. Often by a visual inspection of the velocity field, it was possible to decide whether reasonably good fits were obtained or not. It was found in some cases that the dynamical centre was not coincident with the optical centre (e.g., in IC 1953). The centre and the systemic velocity were computed as un-weighted mean of their values in all the rings for which satisfactory solutions could be obtained. Once the centre and the systemic velocity are determined, these values are held fixed for all radii during successive iterations.

4.5.2 Position angle

In this step, the position angle is determined as a function of radius. The inclination angle for each radius is kept fixed and V_{rot} is allowed to vary. The variations in position angle with radii can be inferred from a visual inspection of the velocity field. If a galaxy is warped, the position angle varies gradually starting at certain radii and often becomes constant at large radii. A warp often results in variations in both the inclination and the position angle simultaneously. For galaxies, where no significant change in the position angle or in the inclination or in both were seen, an average value of the position angle was taken.

4.5.3 *Inclination*

In this iteration, the variations in the inclination angle are modeled as a function of radius. The previously determined parameters (centre, systemic velocity, and PA) are kept fixed. Reliable estimates for the inclination can not be made for galaxies having low inclination (usually below 45°), using the tilted ring model (Begeman 1989). In such cases, optical inclination angles are more reliable.

4.5.4 *Rotation velocities*

In the final iteration, all the disk parameters obtained from the previous iterations are kept fixed, and a fit is carried out to obtain the circular rotation velocity as a function of radius. The receding and the approaching sides of the galaxies were fitted separately. In this step, a cosine weighting scheme was adopted such that points along the major axis have maximum weight and points along the minor axis have zero weight. In addition, points within ± 20 degree from the minor axis were excluded. This is due to the fact that the minor axis does not have any information on the rotation velocity, and the points along the major axis give a direct estimate of the projected rotation velocity.

The rotation velocities obtained from the tilted ring fit should be corrected for the effects of beam smearing. The effects of beam smearing are maximum at the locations of steep velocity gradients. The flat portion of the rotation curve remains nearly unaffected due to beam smearing. The rotation curves presented here are not corrected for the effects of beam smearing. Therefore, for those analyses which are critically dependent on the quality of the rotation curves (e.g., mass modeling), these rotation curves should not be used without making an appropriate correction. Such analyses will be presented elsewhere.

4.6 *Radial H_I density profiles*

The total H_I maps were used to estimate the mean radial H_I surface density profiles by azimuthally averaging the H_I column densities in concentric elliptical annuli. The axial parameters of the ellipses were obtained from the tilted ring models. The elliptical annuli were sampled at intervals of $10''$ to get two points in one synthesised beam of $20''$. The widths of the annuli were kept fixed at $10''$. The average radial profile in each annulus is scaled by the ratio of the total H_I mass to the summed column density in all annuli to obtain the H_I mass surface density. The profiles are corrected for the projection effects to obtain the face-on mass surface density in units of $M_\odot \text{ pc}^{-2}$. The profiles can be quite uncertain in high inclination galaxies where some flux density from the lower radii along the minor axis will be included at larger radii due to finite spatial resolution. Also, the profiles will artificially extend to larger radii near the outer edge of the H_I disk due to the finite angular resolution.

4.7 *Dynamical mass*

The dynamical mass can be estimated from the rotation curves. The total mass within a radius R (kpc) can be derived using the relation $M_{\text{tot}}(M_\odot) = 2.3 \times 10^5 V_{\text{rot}}^2 R / G$,

where V_{rot} is the rotation velocity in km s^{-1} . There can be several choices for the radius R , e.g., $H\text{I}$ disk radius, optical disk radius, disk scale length, etc. The dynamical masses were estimated within the optical radius (i.e., $D_{25}/2$) in the present analysis. The estimates were made only for those galaxies in which flat rotation curves were detected.

5. $H\text{I}$ properties of galaxies

In this section, some of the $H\text{I}$ properties of galaxies derived from the GMRT observations are presented and are compared with those for nearby galaxies in field and in other loose groups.

5.1 $H\text{I}$ and total dynamical mass

The histogram of $H\text{I}$ masses of galaxies detected by GMRT is shown in Fig. 11. The lowest $H\text{I}$ mass ($\sim 8 \times 10^8 M_{\odot}$) detected is that of an S0 galaxy NGC 1481. Some galaxies towards the high mass end in the histogram will have their $H\text{I}$ masses underestimated by GMRT. Due to the limited sample size, this plot is of limited statistical significance to estimate the $H\text{I}$ mass function.

The dynamical masses are plotted in Fig. 12 as a function of Hubble type. Only those galaxies are plotted whose flat part of the rotation curve could be measured. There seems to be a systematic trend in the sense that early type galaxies have on average higher dynamical masses than the late type galaxies. This result is consistent with other studies (e.g., Broeils & Rhee 1997; Verheijen & Sancisi 2001).

The statistical significance of the correlation was estimated using the Spearman Rank-Order Correlation Coefficient method. The significance of the non-zero correlation coefficient in Fig. 12 is $\sim 97\%$.

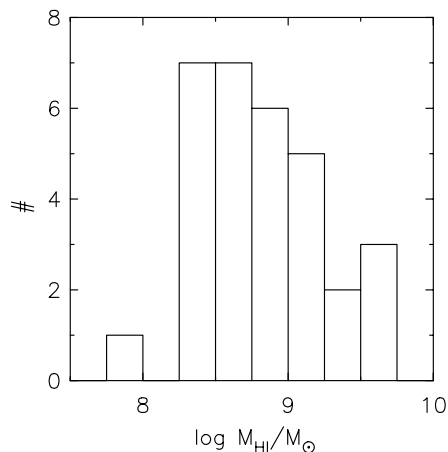


Figure 11. The histogram of $H\text{I}$ masses of the Eridanus galaxies observed with the GMRT.

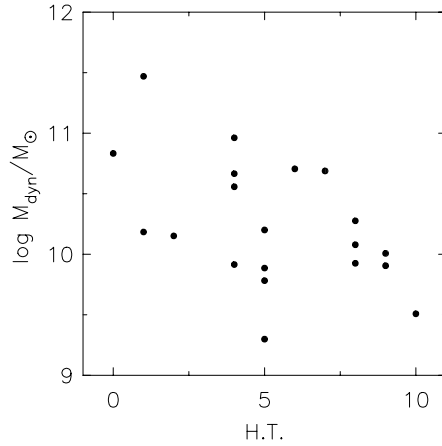


Figure 12. Dynamical mass plotted against the Hubble type. The masses are estimated within the optical diameter (D_{25}).

5.2 H_I mass to luminosity ratio

The ratio of the H_I mass to the luminosities in the B, R, J, and K bands are plotted in the different panels of Fig. 13 respectively against the Hubble types. The B-band luminosities are estimated using the face-on magnitudes from the RC3 catalog, and assuming the Sun's absolute magnitude in the B-band to be 5.48. The values of M_{H_I} / L_B are in general consistent with that of Broeils & Rhee (1997). No significant trend in the M_{H_I} / L ratio with respect to the type is seen in the B-band. The significance of any correlation is less than $\sim 70\%$.

However, significantly higher probabilities ($> 97\%$) were found for a correlation between M_{H_I} / L and H.T. in the other bands. The correlations in the R and the near-IR (J & K) bands are in the sense that late type galaxies have on an average higher M_{H_I} / L value. This trend in M_{H_I} / L in the J and K bands for which the extinction corrections are relatively less significant are consistent with the results for the Ursa-Major galaxies (Verheijen & Sancisi 2001).

5.3 Ratio of H_I diameter to optical diameter

The ratio of the H_I diameter to the optical diameter is plotted in Fig. 14 against various properties of the galaxies. No significant correlation is seen with the Hubble type, the H_I mass, the projected distance to the nearest neighbour, and the J-band magnitude. Contrary to this, Verheijen & Sancisi (2001) found a significant correlation between D_{H_I} / D_{25} and J-band magnitude as well as Hubble type for the Ursa-Major group of galaxies. The average value of D_{H_I} / D_{25} is 1.7 ± 0.8 , which is consistent with the value of 1.7 ± 0.5 obtained by Broeils & Rhee (1997) for a sample of nearby luminous galaxies.

5.4 H_I surface density

The H_I masses of galaxies are observed to correlate with their H_I diameters and optical diameters (Haynes & Giovanelli 1984; Broeils & Rhee 1997; Verheijen & Sancisi

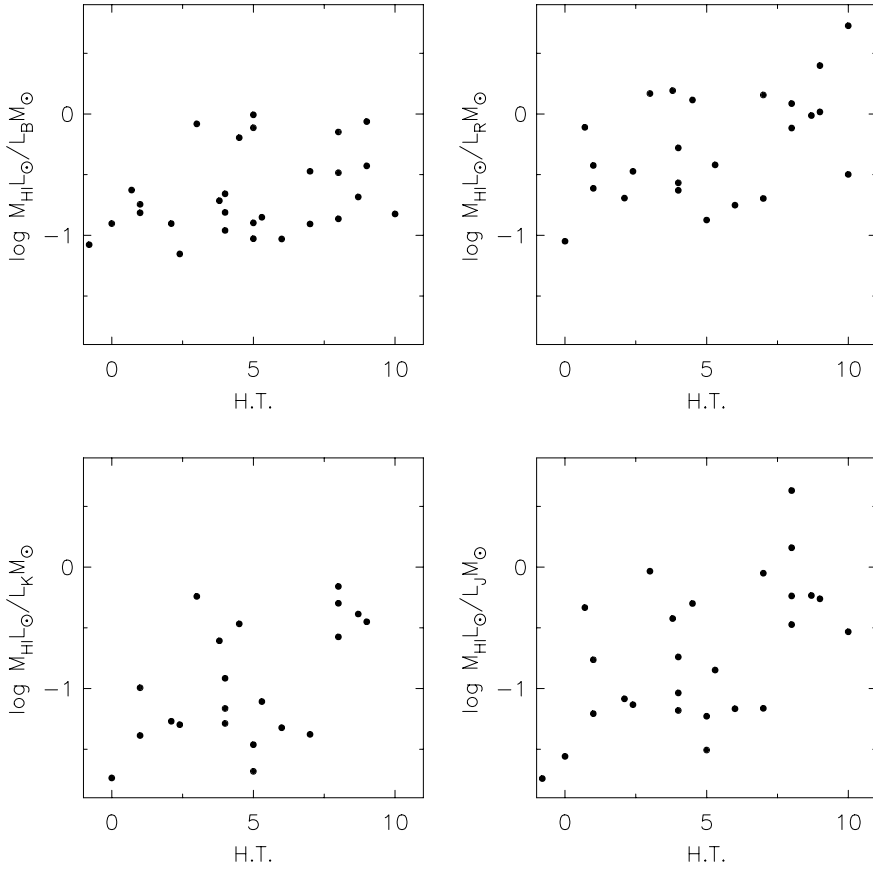


Figure 13. The H_{I} mass to luminosity ratio in the optical and the near-IR bands. There appears to be a trend in the R, and in the near-IR bands in the sense that late type galaxies have on an average higher value of $M_{H_{\text{I}}} / L$.

2001). The latter authors showed that the correlation is tighter with the H_{I} diameter than with the optical diameter with a slope of ~ 2 . The lower panel of Fig. 15 shows $\log M_{H_{\text{I}}}$ for the Eridanus galaxies plotted against $\log D$ for both optical and H_{I} diameters. A straight line is fitted with a fixed slope of 2. It can be seen that the H_{I} diameters of the galaxies in the Eridanus group are more tightly correlated than the optical diameters with their H_{I} masses. The upper panel of Fig. 15 shows the distribution of the H_{I} mass surface densities ($\sigma_{H_{\text{I}}} = 4M_{H_{\text{I}}} / \pi D^2$) of Eridanus galaxies against their Hubble types. It can be seen that the mass surface densities over the H_{I} disks are less scattered compared to the mass surface density over the optical disks.

A summary of the various estimated parameters of the Eridanus galaxies can be found in Table 5.

6. Summary

The Eridanus group of galaxies were observed in the H_{I} 21 cm-line using the GMRT for ~ 200 h. There is considerable scope to use this data for several studies which will

be presented in subsequent papers. The following conclusions can be drawn from the H I properties of the Eridanus galaxies:

- The early type disk galaxies seem to have higher dynamical masses compared to the late type disk galaxies.
- The $M_{\text{H I}}/L$ ratio shows a trend with respect to the type in the R-band and in the near-IR (J & K) bands in the sense that late type galaxies have, on an average, higher $M_{\text{H I}}/L$.
- The average value of $D_{\text{H I}}/D_{\text{opt}}$ is 1.7 ± 0.8 for the Eridanus galaxies is consistent with that for galaxies in other groups and fields.
- The ratio of $M_{\text{H I}}/D_{\text{H I}}$ has less scatter compared to the ratio of $M_{\text{H I}}/D_{\text{opt}}$ for different galaxy types.

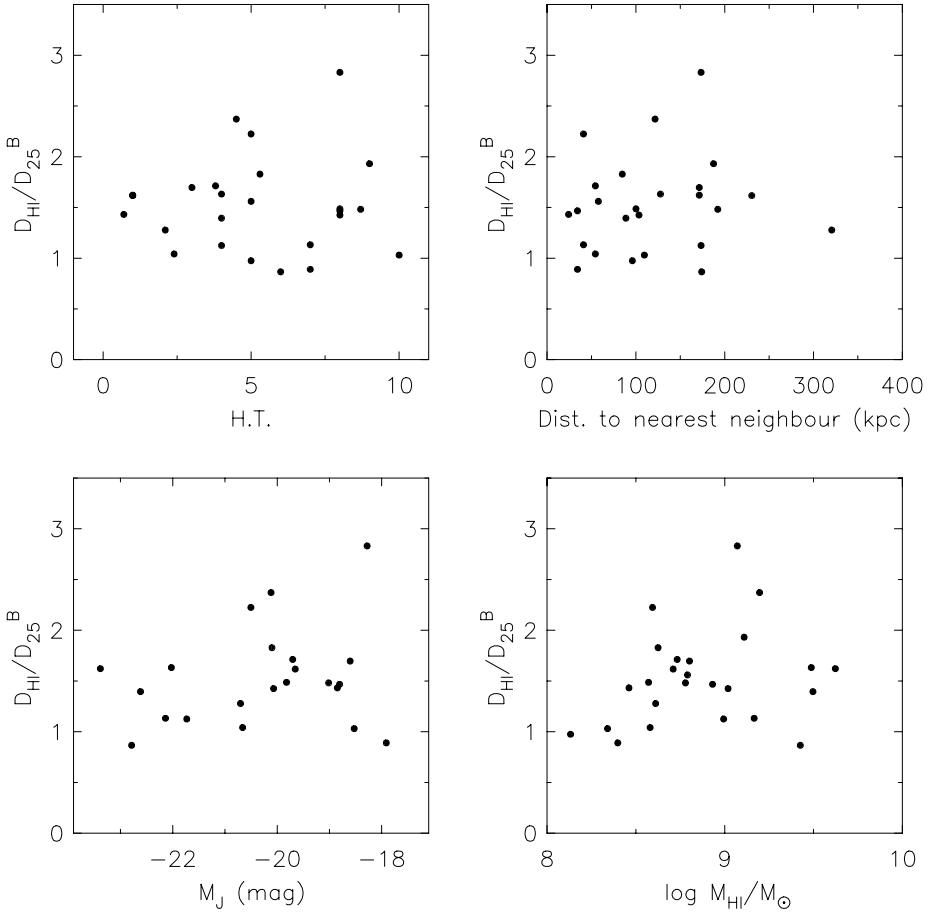


Figure 14. The ratio of the H I diameter to the optical diameter is plotted with different properties of galaxies. The mean ratio of the sample is 1.7 ± 0.8 . This ratio is independent of the Hubble type, the H I mass, the distance to the nearest neighbour, and the J-band luminosity.

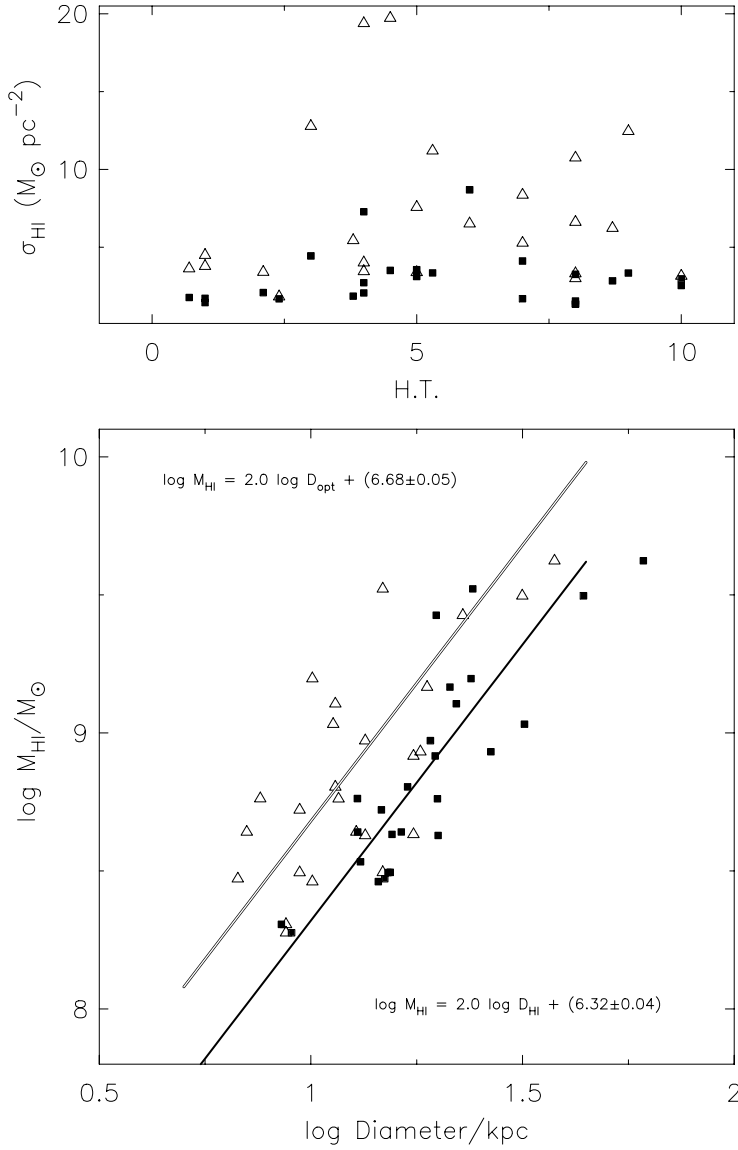


Figure 15. (Lower panel) The H I mass is plotted against H I diameter (filled squares) and optical diameter (open triangles). The two lines are the best fits to the two sets of the data respectively. The slopes were kept fixed at 2. (Upper panel) The H I mass surface density estimated over the H I disk (filled squares) and over the optical disk (open triangles) plotted against the Hubble type.

Acknowledgements

AO would like to acknowledge the encouragement and the support of late K. R. Anantharamaiah in beginning this observational program with the GMRT. We thank Marc Verheijen, Jayaram Chengalur, and the referee for useful comments on various issues discussed here. We thank the staff of the GMRT who made these

observations possible. The GMRT is operated by the National Centre for Radio Astrophysics of the Tata Institute of Fundamental Research. This research has been benefited by the NASA's Astrophysics Data System (ADS) and Extra-galactic Database (NED) services. Their open software used in this research is greatly acknowledged.

Appendix A: The Eridanus group of galaxies – optical, infrared and H_I properties

Table 2. Sample of galaxies.

#	Name	h	α (J2000) m s	δ (J2000) ° ' "	l_{SG} °	b_{SG} °	cz (km/s)
1	NGC 1297	03	19 14.2	−19 06 00	284.30	−38.61	1579
2	NGC 1309	03	22 06.5	−15 24 00	289.12	−38.86	2135
3	NGC 1315	03	23 06.6	−21 22 31	281.46	−39.68	1673
4	SGC 0321.2-1929	03	23 25.1	−19 17 00	284.18	−39.61	1545
5	UGCA 068	03	23 47.2	−19 45 15	283.58	−39.73	1838
6	NGC 1325	03	24 25.4	−21 32 36	281.26	−39.99	1589
7	ESO 548-G 016	03	26 02.4	−21 20 26	281.55	−40.36	2119
8	NGC 1332	03	26 17.3	−21 20 07	281.56	−40.41	1524
9	NGC 1331	03	26 28.3	−21 21 20	281.54	−40.46	1210
10	APMBGC						
	548 + 070 + 070	03	26 31.3	−21 13 01	281.72	−40.46	1548
11	ESO 548-G 021	03	27 35.3	−21 13 42	281.72	−40.71	1668
12	ESO 548-G 025	03	29 00.7	−22 08 45	280.53	−41.08	1680
13	NGC 1345	03	29 31.7	−17 46 40	286.32	−40.91	1529
14	NGC 1347	03	29 41.8	−22 16 45	280.35	−41.24	1759
15	NGC 1353	03	32 03.0	−20 49 09	282.34	−41.73	1525
16	UGCA 077	03	32 19.2	−17 43 05	286.48	−41.57	1961
17	ESO 482-G 005	03	33 02.2	−24 07 58	277.89	−42.03	1915
18	IC 1952	03	33 26.7	−23 42 46	278.46	−42.13	1812
19	ESO 548-G 036	03	33 27.6	−21 33 53	281.35	−42.09	1480
20	IC 1953	03	33 41.9	−21 28 43	281.47	−42.14	1867
21	NGC 1359	03	33 47.7	−19 29 31	284.15	−42.06	1966
22	ESO 548-G 043	03	34 10.5	−19 33 30	284.07	−42.16	1931
23	ESO 548-G 044	03	34 19.2	−19 25 28	284.25	−42.18	1696
24	NGC 1371	03	35 02.0	−24 55 59	276.80	−42.48	1471
25	NGC 1370	03	35 14.6	−20 22 25	282.99	−42.46	1063
26	ESO 548-G 049	03	35 28.1	−21 13 01	281.85	−42.55	1510

Table 2. (Continued)

#	Name	$\alpha(\text{J2000})$			$\delta(\text{J2000})$			l_{SG}	b_{SG}	cz
		h	m	s	°	'	"	°	°	(km/s)
27	IC 1962	03	35	37.5	−21	17	39	281.74	−42.58	1806
28	NGC 1377	03	36	39.1	−20	54	08	282.29	−42.81	1792
29	ESO 482-G 013	03	36	53.9	−24	54	46	276.82	−42.90	1835
30	NGC 1385	03	37	28.3	−24	30	05	277.38	−43.04	1493
31	NGC 1383	03	37	39.2	−18	20	22	285.80	−42.89	1948
32	NGC 1390	03	37	52.2	−19	00	30	284.89	−42.99	1207
33	NGC 1393	03	38	38.6	−18	25	41	285.70	−43.13	2185
34	NGC 1401	03	39	21.8	−22	43	29	279.81	−43.48	1495
35	ESO 548-G 064	03	40	00.0	−19	25	35	284.36	−43.52	1694
36	ESO 548-G 065	03	40	02.7	−19	22	00	284.45	−43.53	1221
37	IC 0343	03	40	07.1	−18	26	36	285.72	−43.48	1841
38	NGC 1407	03	40	11.9	−18	34	49	285.53	−43.51	1779
39	ESO 482-G 031	03	40	41.5	−22	39	04	279.92	−43.79	1621
40	APMBGC									
	548-110-078	03	40	52.5	−18	28	29	285.69	−43.66	1595
41	NGC 1415	03	40	56.7	−22	33	47	280.04	−43.85	1585
42	NGC 1414	03	40	57.0	−21	42	47	281.22	−43.84	1681
43	ESO 548-G 072	03	41	00.3	−19	27	19	284.34	−43.76	2034
44	NGC 1416	03	41	02.9	−22	43	09	279.82	−43.87	2167
45	IC 0345	03	41	09.1	−18	18	51	285.92	−43.71	1335
46	ESO 482-G 035	03	41	15.0	−23	50	10	278.27	−43.91	1890
47	NGC 1422	03	41	31.1	−21	40	54	281.27	−43.97	1637
48	IC 0346	03	41	44.6	−18	16	01	286.00	−43.85	2013
49	ESO 549-G 002	03	42	57.3	−19	01	12	284.99	−44.19	1111
50	MCG-03-10-041	03	43	35.5	−16	00	52	289.18	−44.07	1215
51	NGC 1440	03	45	02.9	−18	15	58	286.09	−44.63	1534
52	NGC 1438	03	45	17.2	−23	00	09	279.43	−44.85	1555
53	NGC 1452	03	45	22.3	−18	38	01	285.58	−44.73	1737
54	ESO 549-G 018	03	48	14.1	−21	28	28	281.60	−45.52	1587
55	NGC 1481	03	54	28.9	−20	25	38	283.15	−46.96	1733
56	NGC 1482	03	54	39.3	−20	30	09	283.04	−47.00	1916
57	ESO 549-G 035	03	55	04.0	−20	23	01	283.22	−47.10	1778

Note: The positions and velocities are from the NASA Extra-galactic Database (NED).

Column 1: Serial number; **Column 2:** Name of the galaxy; **Columns 3 & 4:** Equatorial coordinates in the J2000 epoch; **Columns 5 & 6:** Super Galactic coordinates; **Column 7:** Heliocentric velocity (optical definition).

Table 3. Optical and H_I properties of galaxies.

#	Name	Morph.	H.T.	B_0^T (mag)	K^T (mag)	B – K (mag)	H _I Flux (Jy km/s)	W_{20} (km/s)
1	NGC 1297	S0a	−2.3	12.65	8.93	3.72		
2	NGC 1309	Sbc	4.0	11.83	9.10	2.73	18.7	171
3	NGC 1315	S0	−1.0	13.38	9.73	3.65		
4	SGC 0321.2-1929	Im	10.0				3.1	79
5	UGCA 068	Scdm	8.7	13.56			5.9	131
6	NGC 1325	Sbc	4.0	11.51	8.63	2.88	24.4	348
7	ESO 548-G 016	S?		14.61				
8	NGC 1332	S0	−3.0	11.21	7.05	4.16		
9	NGC 1331	E/S0	−5.0	14.22	10.76	3.46		
10	APMBGC 548 + 070 + 070	S0			11.23			
11	ESO 548-G 021	Sdm						
12	ESO 548-G 025	Sa	0.7	14.47				
13	NGC 1345	Sc	4.5	13.80			12.0	138
14	NGC 1347	Scd	5.3	13.55			0.6	105
15	NGC 1353	Sbc	3.0	11.73	8.11	3.62		
16	UGCA 077	Sdm	9.0	14.36			5.3	158
17	ESO 482-G 005	Sdm	8.0	14.33			5.4	184
18	IC 1952	Sbc	4.0	12.59	9.87	2.72	5.1	263
19	ESO 548-G 036	S?			10.43		4.1	
20	IC 1953	Sc	7.0	12.10	9.65	2.45	8.1	197
21	NGC 1359	Scm	9.0	12.37	11.17	1.20	25.7	223
22	ESO 548-G 043	Sa		15.55	11.46	4.09		
23	ESO 548-G 044	S0/a	−1.3	14.11	10.37	3.74		
24	NGC 1371	Sa	1.0	11.36	7.63	3.73	53.7	427
25	NGC 1370	E/S0	−3.5	13.36	9.87	3.49		
26	ESO 548-G 049	S?		14.92				
27	IC 1962	Sdm	8.0	13.74			5.5	171
28	NGC 1377	S0	−2.0	13.29	9.72	3.57		
29	ESO 482-G 013	Sb					2.0	99
30	NGC 1385	Scd	6.0	11.14	8.57	2.57	22.6	204
31	NGC 1383	S0	−2.0	13.25	9.44	3.81		
32	NGC 1390	S0/a	1.0	14.01	11.54	2.47	2.4	191
33	NGC 1393	S0	−1.9	12.78	9.18	3.60		

Table 3. (Continued)

#	Name	Morph.	H.T.	B_0^T (mag)	K^T (mag)	B – K (mag)	H I Flux (Jy km/s)	W_{20} (km/s)
34	NGC 1401	S0	–2.0	13.11	9.35	3.76		
35	ESO 548-G 064	S0		14.52	10.73	3.79		
36	ESO 548-G 065	Sa	0.7	14.56	12.90	1.66	2.0	145
37	IC 0343	S0	–1.0	13.91	10.50	3.41		
38	NGC 1407	E0	–5.0	10.71	6.70	4.01		
39	ESO 482-G 031	dS0			12.37			
40	APMBGC							
	548-110-078	dS0						
41	NGC 1415	S0/a	0.0	12.41	8.73	3.68		
42	NGC 1414	Sbc	3.8	13.59			2.1	
43	ESO 548-G 072	S?						
44	NGC 1416	E/S0	–5.0	13.88	10.54	3.34		
45	IC 0345	S0/a		14.64	10.05	4.59		
46	ESO 482-G 035	Sab	2.1	13.42			4.2	177
47	NGC 1422	Sab	2.4	13.16	10.73	2.43		
48	IC 0346	S0	–0.8	13.37	9.78	3.59		
49	ESO 549-G 002	Im	10.0	14.53				
50	MCG-03-10-041	Sdm	8.0		11.91		3.7	184
51	NGC 1440	S0	–1.9	12.35	8.20	4.15		
52	NGC 1438	S0/a	0.0	12.94	9.62	3.32		
53	NGC 1452	Sa	0.4	12.56	8.67	3.89		
54	ESO 549-G 018	Sc	5.0	13.13	10.64	2.49		
55	NGC 1481	S0	–3.3	14.40	11.18	3.22		
56	NGC 1482	S0/a	–0.8	13.01	8.48	4.53	5.5	131
57	ESO 549-G 035	Sc					6.0	145

Note: The morphological types, Hubble types, and photographic B-band magnitudes are from RC3 (third reference catalog of galaxies; de Vaucouleurs *et al.* 1991) provided by NED. The K-band magnitudes are from the *Two Micron All Sky Survey* (2MASS; Jarrett *et al.* 2000). The single dish H I flux integrals and H I widths are from NED.

Column 1: Serial number; **Column 2:** Name of the galaxy; **Column 3:** Morphological type; **Column 4:** Hubble type; **Column 5:** Apparent total photographic B-band magnitude corrected for inclination; **Column 6:** Apparent total K-band magnitude from 2MASS; **Column 7:** B–K color; **Columns 8 & 9:** Single dish H I flux integral (Jy km s^{–1}) and velocity width (km s^{–1}) at 20% of the peak flux density in the global H I profile.

Table 4. Observational details.

#	Date (dd-mm-yy)	Pointing centre		Galaxy name	Seps. (arc min)	rms (mJy/ beam)
		α (J2000) (h m s)	δ (J2000) ($^{\circ}$ ' ")			
1	01-03-00	03 33 47.7	-19 29 31.0	N 1359	0	1.8
2	27-10-00	03 35 01.4	-24 55 58.0	N 1371	0	1.3
3	27-10-00	03 37 28.3	-24 30 05.0	N 1385	0	1.3
4	15-12-00	03 22 06.5	-15 24 00.0	N 1309	0	1.7
5	03-05-01	03 24 25.6	-21 32 39.0	N 1325	0	1.9
6	06-05-01	03 23 25.1	-19 17 04.0	S 0321.2-1929	0	1.3
7	17-05-01	03 54 33.4	-20 28 14.2	N 1482	2.3	1.0
				N 1481	2.8	1.1
				E 549-G 035	8.9	1.4
8	18-05-01	03 40 34.4	-18 30 23.0	A 548-110-078 ($\times$)	4.6	1.0
				N 1407 ($\times$)	6.9	1.1
				I 0343 ($\times$)	7.5	1.1
				A 548-108-069 ($\times$)	8.6	1.1
9	19-05-01	03 27 02.3	-21 16 05.7	N 1332 ($\times$)	4.1	1.1
				E 548-G 016 ($\times$)	5.6	1.1
				N 1331 ($\times$)	5.9	1.1
10	20-05-01	03 45 02.9	-18 15 58.0	N 1440 ($\times$)	0	0.9
11	21-05-01	03 35 14.6	-20 22 25.0	N 1370 ($\times$)	0	0.9
12	22-05-01	03 36 39.1	-20 54 08.0	N 1377 ($\times$)	0	0.9
13	16-06-02	03 29 31.7	-17 46 43.0	N 1345	0	1.5
14	16-06-02	03 32 19.3	-17 43 07.0	U 077	0	1.4
15	17-06-02	03 35 37.3	-21 17 39.1	I 1962	0	0.9
				E 548-G 049	5.1	1.1
16	17-06-02	03 23 47.2	-19 45 15.0	U 068	0	0.9
17	18-06-02	03 41 13.9	-21 41 48.5	N 1414	4.0	0.9
				N 1422	4.1	1.0
18	20-06-02	03 40 56.8	-22 38 24.5	E 482-G 031 ($\times$)	3.6	0.9
				N 1415	4.5	0.9
				N 1416 ($\times$)	4.9	0.9
19	22-06-02	03 29 21.1	-22 12 43.9	E 548-G 025 ($\times$)	6.2	1.2
				N 1347	6.3	1.2
20	23-06-02	03 33 26.5	-23 42 41.0	I 1952	0	1.2
21	23-06-02	03 33 34.5	-21 30 58.7	I 1953	2.8	1.4
				E 548-G 036	3.3	1.4

Table 4. (*Continued*)

#	Date (dd-mm-yy)	Pointing centre		Galaxy name	Sepa. (arc min)	rms (mJy/ beam)
		α (J2000) (h m s)	δ (J2000) ($^{\circ}$ ' ")			
22	24-06-02	03 27 35.3	−21 13 42.0	E 548-G 021	0	0.9
23	24-06-02	03 37 52.2	−19 00 30.0	N 1390	0	1.1
24	26-06-02	03 41 15.0	−23 50 10.0	E 482-G 035	0	1.2
25	26-06-02	03 48 14.1	−21 28 28.0	E 549-G 018	0	1.4
26	27-06-02	03 40 31.4	−19 24 40.0	E 548-G 065	7.3	1.1
				E 548-G 072	7.3	1.3
				E 548-G 064 (×)	7.5	1.3
27	27-06-02	03 42 57.3	−19 01 12.0	E 549-G 002	0	1.0
28	28-06-02	03 41 00.3	−19 27 19.0	E 548-G 077 (×)	0	1.2
29	28-06-02	03 43 35.5	−16 00 52.0	M -03-10-041	0	1.2
30	29-06-02	03 33 02.2	−24 07 58.0	E 482-G 005	0	1.1
31	29-06-02	03 36 53.9	−24 54 46.0	E 482-G 013	0	1.1

Notes: (1) The bandwidth was 8 MHz except for pointings 5 and 6 where it was 4 MHz and for pointing 4 it was 2 MHz. Therefore, the velocity resolutions were $\sim 13.4 \text{ km s}^{-1}$ for all except for N 1325 and S 0321.2-1929 where it was $\sim 6.7 \text{ km s}^{-1}$ and for N 1309 it was $\sim 3.3 \text{ km s}^{-1}$.

(2) The on-source integration time for observation was $\sim 3 \text{ h}$ except for pointings 5, 6, 7, 8, 9 where it was $\sim 7 \text{ h}$.

(3) The centre frequency was 1412.6 MHz except for pointing 1 (1408.0 MHz), pointings 2 and 3 (1414.0 MHz), pointing 4 (1410.0 MHz), pointings 5 and 6 ($\sim 1413 \text{ MHz}$) and pointing 7 (1412.0 MHz).

(4) The column with the heading ‘Sepa’ indicates the angular separation between the pointing centre and the galaxy.

(5) The *FWHM* of the image cubes was $20'' \times 20''$ except for N 1309 and N 1325 where it was $25'' \times 25''$ and for N 1371, E 482-G 031, N 1415, and N 1416 it was $30'' \times 30''$.

(6) A cross against the galaxy name indicates an H I non-detection.

Column 1: Serial number; **Column 2:** Date of observation; **Column 3:** Coordinates of the pointing centre; **Column 4:** Galaxies in the field of view; **Column 5:** Separation (arc min) of galaxy(ies) in the field of view from the pointing centre; **Column 6:** RMS noise (mJy beam $^{-1}$) in the channel images.

Table 5. Results from the GMRT HI synthesis observations.

Galaxy name							
Opt. centre		V_{sys}	D_{HI}	V_{max}	M_{HI}	M_{HI} / L_B	$\sigma_{\text{HI}} (\text{HI})$
Radio centre	Type	V_{cent}	D_{B25}	V_{flat}	M_K	M_{HI} / L_R	$\sigma_{\text{HI}} (\text{opt})$
$(\alpha(h, m, s),$	INCL	V_{opt}	D_{R25}	$W_{50}/2$	M_{dyn}	M_{HI} / L_K	$D_{\text{HI}} / D_{\text{B25}}$
$\delta(^{\circ}, ', ') \text{ J2000})$	PA	(km/s)	(kpc)	(km/s)	($10^9 M_{\odot}$)	($M_{\odot}/L_{\odot}$)	($M_{\odot}/\text{pc}^2$)
ESO 482-G 005	SBdm	1923	32	80.2	1.08	0.71	1.34
03 33 02.2 -24 07 58	82°	1918	11.3	80.2	–	–	10.7
03 33 2.15 -24 07 58	264°	1915	–	79.5	8.423	–	2.83
ESO 482-G 013	Sb	–	12.9	–	0.578	0.83	4.44
03 36 53.9 -24 54 46	63	1850	7.59	–	0.604	1.5	12.8
03 36 53.8 -24 54 46	65	1835	7.17	–	1.11	0.57	1.7
ESO 482-G 035	SBab	1883	16.4	118	0.438	0.13	2.08
03 41 15.0 -23 50 10	49	1884	12.8	118	4.883	0.2	3.4
03 41 14.7 -23 50 11	185	1890	14.8	115	20.78	0.054	1.28
ESO 548-G 021	SBdm	1690	20	87.7	0.425	0.14	1.36
03 27 35.3 -21 13 42	80	1691	13.4	87.7	0.507	–	3
03 27 35.2 -21 13 41.7	64	1668	–	83.3	11.99	0.5	1.49
ESO 548-G 036	S?	–	–	–	0.257	0.13	–
03 33 27.6 -21 33 53	–	1507	6.7	–	4.482	–	7.31
03 33 28 -21 33 55.1	–	1480	–	60.5	–	0.034	–
ESO 548-G 049	S?	1533	14.9	71	0.296	0.34	1.69
03 35 28.1 -21 13 01	71	1533	6.72	71	–	1.4	8.36
03 35 28.4 -21 13 7.01	128	1510	6.27	–	3.93	–	2.22
ESO 548-G 065	SBa	1213	14.4	72.9	0.289	0.24	1.77
03 40 02.7 -19 22 00	80	1243	10.1	72.9	–	0.78	3.63
03 40 2.53 -19 21 56.8	37	1221	–	64.5	6.214	–	1.43
ESO 548-G 072	S?	2045	8.52	44.3	0.203	0.77	3.56
03 41 00.3 -19 27 19	74	2052	8.74	44.3	–	–	3.38
03 41 0.795 -19 27 19	51	2034	–	48.0	1.989	–	0.975

Table 5. (Continued)

Galaxy name							
Opt. centre		V_{sys}	$D_{\text{H I}}$	V_{max}	$M_{\text{H I}}$	$M_{\text{H I}}/L_B$	$\sigma_{\text{H I}}(\text{H I})$
Radio centre	Type	V_{cent}	D_{B25}	V_{flat}	M_K	$M_{\text{H I}}/L_R$	$\sigma_{\text{H I}}(\text{opt})$
$(\alpha(h, m, s),$	INCL	V_{opt}	D_{R25}	$W_{50}/2$	M_{dyn}	$M_{\text{H I}}/L_K$	$D_{\text{H I}}/D_{\text{B25}}$
$\delta(^{\circ}, ', ") \text{ J2000})$	PA	(km/s)	(kpc)	(km/s)	($10^9 M_{\odot}$)	($M_{\odot}/L_{\odot}$)	($M_{\odot}/\text{pc}^2$)
ESO 549-G 002	IBm	1110	9	55.9	0.189	0.15	2.97
03 42 57.3 -19 01 12	63	1115	8.74	55.9	–	0.32	3.15
03 42 57.2 -19 01 11.1	210	1111	–	60.8	3.167	–	1.03
ESO 549-G 018	SABc	1587	15.5	–	0.429	0.094	2.26
03 48 14.1 -21 28 28	56	1576	17.5	–	12.37	0.13	1.79
03 48 14 -21 28 28.9	203	1587	17.5	152	–	0.021	0.89
ESO 549-G 035	Sc	1814	14.7	74.5	0.526	0.98	3.11
03 55 04.0-20 23 01	56	1768	9.41	74.5	–	–	7.57
03 55 4.39-20 23 0.92	30	1778	–	81.3	6.057	–	1.56
IC 1952	SBbc	1820	19.7	134	0.825	0.11	2.72
03 33 26.7 -23 42 46	71	1823	17.5	134	9.599	0.23	3.44
03 33 26.4 -23 42 46.1	319	1812	22.4	122	36.12	0.052	1.13
IC 1953	SBd	1863	21.3	150	1.46	0.12	4.11
03 33 41.9 -21 28 43	37	1863	18.8	150	20.97	0.2	5.27
03 33 42.2 -21 28 39.3	129	1867	24.2	142	48.78	0.042	1.13
IC 1962	SBdm	1806	26.6	82.3	0.855	0.33	1.54
03 35 37.4 -21 17 33	80	1811	18.1	82.3	0.7401	1.2	3.31
03 35 37.8 -21 17 38	358	1806	18.4	69.1	14.26	0.69	1.47
MCG-03-10-041	SBdm	1207	19.2	110	0.938	–	3.25
03 43 35.5 -16 00 52	57	1217	13.4	110	2.112	0.77	6.61
03 43 35.4 -16 0 51.3	343	1215	13.9	109	18.9	0.27	1.43
NGC 1309	SABc	2134	24.1	164	3.33	0.22	7.28
03 22 06.5 -15 24 00	20	2134	14.8	164	16.44	0.53	19.4
03 22 5.89 -15 23 59.9	210	2135	13.4	168	46.35	0.12	1.63

Table 5. (*Continued*)

Galaxy name Opt. centre Radio centre ($\alpha(h, m, s)$, $\delta(^{\circ}, ', ")$ J2000)	Type INCL PA	V_{sys} V_{cent} V_{opt} (km/s)	D_{H1} D_{B25} D_{R25} (kpc)	V_{max} V_{flat} $W_{50}/2$ (km/s)	M_{H1} M_{K} M_{dyn} ($10^9 M_{\odot}$)	$M_{\text{H1}}/L_{\text{B}}$ $M_{\text{H1}}/L_{\text{R}}$ $M_{\text{H1}}/L_{\text{K}}$ ($M_{\odot}/L_{\odot}$)	$\sigma_{\text{H1}}(\text{H1})$ $\sigma_{\text{H1}}(\text{opt})$ $D_{\text{H1}}/D_{\text{B25}}$ ($M_{\odot}/\text{pc}^2$)
NGC 1325	SABc	1593	44	158	3.14	0.15	2.06
03 24 25.6 -21 32 39	71	1594	31.6	158	27.49	0.27	4.01
03 24 25.6 -21 32 35.9	232	1589	38.5	169	91.69	0.069	1.39
NGC 1345	SBc	1530	23.9	116	1.57	0.64	3.51
03 29 31.7 -17 46 43	34	1536	10.1	116	2.766	1.3	19.7
03 29 31.8 -17 46 40.3	88	1529	8.51	112	15.87	0.34	2.37
NGC 1347	SBcd	1758	12.9	96.9	0.438	0.14	3.35
03 29 41.8 -22 16 45	26	1784	7.06	96.9	3.364	0.38	11.2
03 29 41.7 -22 16 44.7	328	1759	8.51	—	7.685	0.078	1.83
NGC 1359	SBcm	1973	—	—	3.44	0.37	—
03 33 47.7 -19 29 31	53	1980	—	—	5.816	1	—
03 33 47.2 -19 29 19.1	325	1966	17.9	108	—	0.35	—
NGC 1371	SABa	1467	61	260	4.2	0.18	1.44
03 35 01.4 -24 55 58	49	1469	37.6	260	61.55	0.24	3.78
03 35 1.89 -24 55 58.6	136	1471	34.5	246	295.1	0.041	1.62
NGC 1385	SBcd	1493	19.8	138	2.67	0.093	8.69
03 37 28.3 -24 30 05	40	1503	22.8	138	33.66	0.18	6.51
03 37 27.8 -24 30 6.68	181	1493	30.5	140	50.77	0.048	0.866
NGC 1390	SB0/a	1207	15.2	118	0.312	0.15	1.72
03 37 52.2 -19 00 30	60	1221	9.41	118	1.845	0.38	4.49
03 37 52.1 -19 0 29.6	24	1207	8.96	97.2	15.27	0.1	1.62

Table 5. (Continued)

Galaxy name							
Opt. centre		V_{sys}	$D_{\text{H I}}$	V_{max}	$M_{\text{H I}}$	$M_{\text{H I}}/L_B$	$\sigma_{\text{H I}} (\text{H I})$
Radio centre	Type	V_{cent}	D_{B25}	V_{flat}	M_K	$M_{\text{H I}}/L_R$	$\sigma_{\text{H I}} (\text{opt})$
$(\alpha(h, m, s),$	INCL	V_{opt}	D_{R25}	$W_{50}/2$	M_{dyn}	$M_{\text{H I}}/L_K$	$D_{\text{H I}}/D_{\text{B25}}$
$\delta(^{\circ}, ', ") \text{ J2000})$	PA	(km/s)	(kpc)	(km/s)	($10^9 M_{\odot}$)	($M_{\odot}/L_{\odot}$)	($M_{\odot}/\text{pc}^2$)
NGC 1414	SBbc	1695	19.9	78.1	0.577	0.19	1.85
03 40 57.0 -21 42 47	80	1660	11.6	78.1	1.4	1.6	5.44
03 40 57.3 -21 42 49.3	357	1681	11.2	95.4	8.226	0.25	1.71
NGC 1415	SAB0/a	1585	–	158	1.11	0.13	–
03 40 56.8 -22 33 52	–	1576	–	158	36.34	0.089	–
03 40 57 -22 33 49.7	–	1585	32.7	167	–	0.018	–
NGC 1422	SBab	1630	15.4	73.8	0.312	0.07	1.68
03 41 31.2 -21 40 53	80	1660	14.8	73.8	3.722	0.34	1.82
03 41 31.2 -21 40 54.4	65	1637	18.4	54.2	9.34	0.05	1.04
NGC 1481	S0	–	3.4	–	0.079	0.06	8.8
03 54 29.0 -20 25 38	–	–	6.7	–	2.2	–	2.3
	–	1733	–	–	–	0.02	0.51
NGC 1482	S0/a	–	9.0	–	0.428	0.07	5.0
03 54 38.9 -20 30 08	–	–	16.8	–	28.0	–	1.45
	–	1916	–	–	–	0.01	0.54
UGCA 068	SABcdm	1852	16.9	87.6	0.638	0.21	2.84
03 23 47.2 -19 45 15	34	1843	11.4	87.6	0.931	0.97	6.22
03 23 47.1 -19 45 15.2	35	1838	10.3	84.1	10.17	0.41	1.48
UGCA 077	SBdm	1963	22.1	77.9	1.28	0.87	3.34
03 32 19.3 -17 43 07	66	1956	11.4	77.9	–	2.5	12.4
03 32 19.1 -17 43 10.2	149	1961	12.1	74.5	8.042	–	1.93

Table 5. (Continued)

Galaxy name							
Opt. centre		V_{sys}	$D_{\text{H I}}$	V_{max}	$M_{\text{H I}}$	$M_{\text{H I}} / L_B$	$\sigma_{\text{H I}} (\text{H I})$
Radio centre	Type	V_{cent}	D_{B25}	V_{flat}	M_K	$M_{\text{H I}} / L_R$	$\sigma_{\text{H I}} (\text{opt})$
$(\alpha(h, m, s),$	INCL	V_{opt}	D_{R25}	$W_{50}/2$	M_{dyn}	$M_{\text{H I}} / L_K$	$D_{\text{H I}} / D_{\text{B25}}$
$\delta(^{\circ}, ', ") \text{ J2000})$	PA	(km/s)	(kpc)	(km/s)	($10^9 M_{\odot}$)	($M_{\odot}/L_{\odot}$)	($M_{\odot}/\text{pc}^2$)
SGC 0321.2-1929	IBm	1552	13.1	56.4	0.341	–	2.53
03 23 25.1 -19 17 04	24	1556	–	56.4	–	5.3	–
03 23 25.2 -19 17 5.33	175	1545	–	54.7	–	–	–

Column 1: Galaxy name, Opt. Centre, Radio Centre: The abbreviated names of galaxies, the optical centre is from the RC3 catalog and the radio centre is estimated from the tilted ring fit; **Column 2:** Type, INCL, PA: The morphological type, the inclination, and the position angle; **Column 3:** Velocities: The systemic velocities (Heliocentric) obtained from the tilted ring fit, from the global profile fit, and from RC3; **Column 4:** Diameters: The H I diameter, the B-band optical diameter at B 25 mag arc sec⁻² from RC3, and the R-band optical diameter at R 25 mag arc sec⁻² from the current observations; **Column 5:** Rotation velocities: The entries in this column are maximum rot. velocity from the rotation curve, velocity at the flat part of the rotation curve, and maximum rot. velocity estimated from the H I global profile corrected for the broadening due to channel width and random and turbulent motions in galaxies; **Column 6:** Masses: The H I mass, the stellar mass from the K-band light assuming M^*/L ratio of 0.6, and the dynamical mass within B25; **Column 7:** Mass to light ratios: The ratios of the H I mass to the B-band, the R-band, and the K-band luminosities; **Column 8:** Mean H I surface densities and the ratio of the H I diameter to the optical diameter: The mean H I surface density ($M_{\odot}/\text{pc}^2$) over the H I diameter and the optical diameter at B25 mag arc sec⁻². The last entry in this column is the ratio of the H I diameter to the optical diameter at 25 mag arc sec⁻².

Appendix B: The Eridanus group of galaxies – The H I atlases

The results are also presented in the form of an atlas. The layout of the atlas is shown in Fig. 16. The high resolution images (20''–30'') are presented in this atlas. The contents of the panels in the atlas are described below:

Total H I image

Contours of total H I image are overlaid upon a gray-scale image of the galaxy from DSS (Digitized Sky Survey). The contours start at $N_{\text{H I}} = 10^{20} \text{ cm}^{-2}$ with increments of $2 \times 10^{20} \text{ cm}^{-2}$. The first contour is white in color and the rest of them are black. The contour lines become thicker (appearing darker) in two steps for higher H I column densities. One step is at $N_{\text{H I}} = 1.2 \times 10^{21} \text{ cm}^{-2}$ till $N_{\text{H I}} = 2 \times 10^{21} \text{ cm}^{-2}$ and the second step is for higher column densities. The synthesised beam is shown at the bottom left hand corner. A horizontal bar of length 5 kpc is drawn near the upper right hand corner.

Velocity field

The velocity field of the galaxy is shown as iso-velocity contours overlaid upon the gray-scale optical DSS image. The white contours indicate velocities in the approaching half and dark contours indicate velocities in the receding half of the galaxy. The contours are drawn in increments of 10 km s^{-1} . The thick dark contour near the centre of the galaxy represents the systemic velocity of the galaxy estimated from the tilted ring model. A horizontal bar of length 5 kpc is drawn near the upper right hand corner.

H I Global profile

The H I global profile is shown as a function of the Heliocentric velocity (optical definition). The systemic velocity of the galaxy estimated from the tilted ring model is written near the top left hand corner of the panel, and is also marked by an arrow along the velocity axis.

H I mass surface density

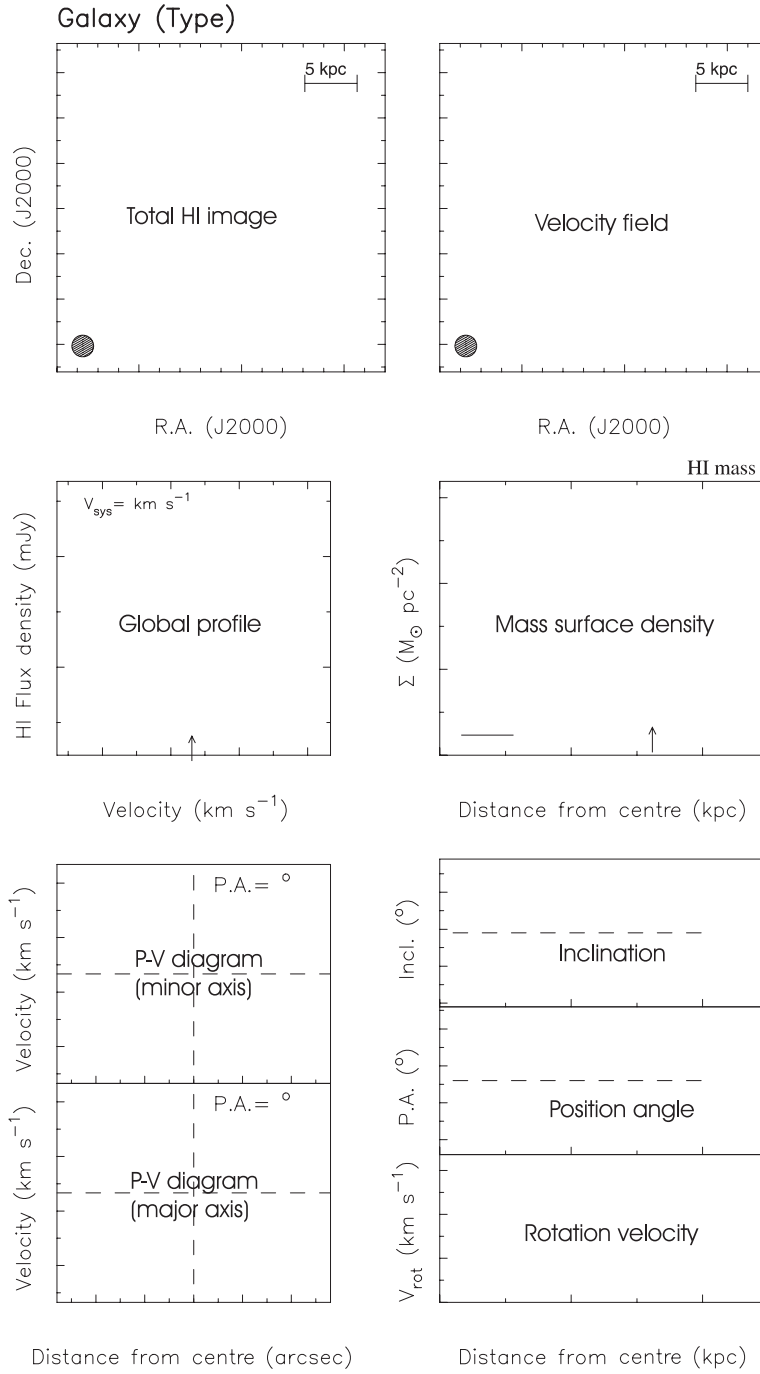
The face-on H I mass surface density is shown in this panel. The values in the approaching and the receding halves are indicated by open and filled circles respectively. The average profile is shown by a thick line. An arrow is marked along the abscissa at the position corresponding to the optical radius estimated at the 25 mag arc sec⁻² B-band isophotal level from the RC3 catalog. A horizontal bar at the left bottom indicates the spatial resolution of the total H I image.

Position–velocity diagrams

The position–velocity (PV) diagrams are shown along the major and the minor axes of the galaxy in the bottom and the top panels respectively. The contours are from 3 mJy beam^{-1} to 8 mJy beam^{-1} in increments of 1 mJy beam^{-1} , and thereafter, in increments of 4 mJy beam^{-1} . A vertical dotted line marks the dynamical centre and a horizontal dotted line marks the systemic velocity, both determined from the tilted ring model. The velocity resolution and the spatial resolution are indicated by the vertical and the horizontal scale-bar respectively at the bottom left hand corners of the two diagrams.

Axial parameters and rotation curve

The axial parameters (PA and INCL) determined from the tilted ring models are plotted in the upper two panels of this figure. The thin lines in these two panels show the final adopted trend of the variations in PA and INCL respectively. The values of the position angle and the inclination determined from the optical images are also marked in the respective panels by an asterisk sign. The rotation curves are shown in the bottom most panel. The approaching and the receding halves are plotted separately using open and filled symbols respectively. The average rotation curve is drawn by a thin line. These rotation curves are not corrected for the effects of beam smearing.

**Figure 16.** The layout of the atlas.

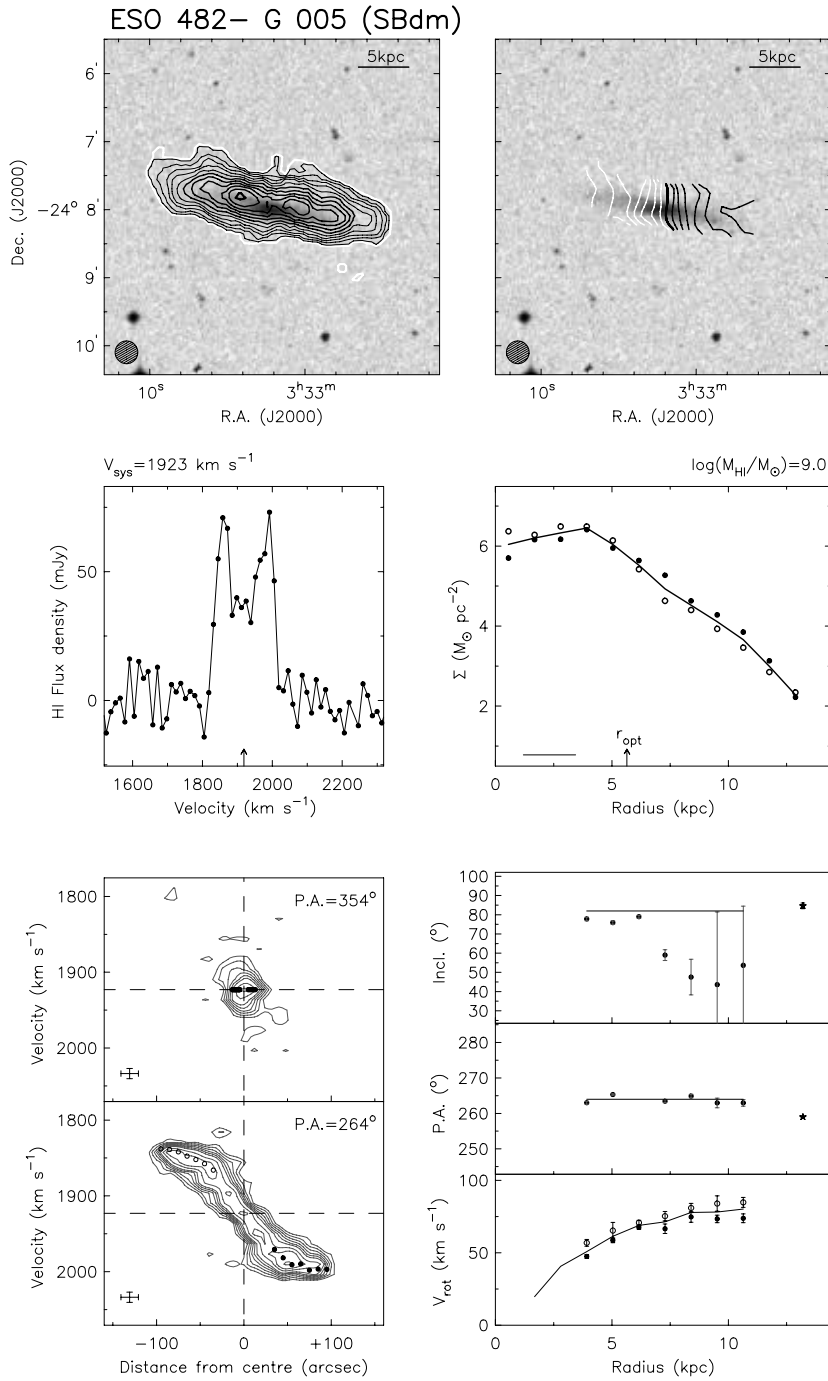
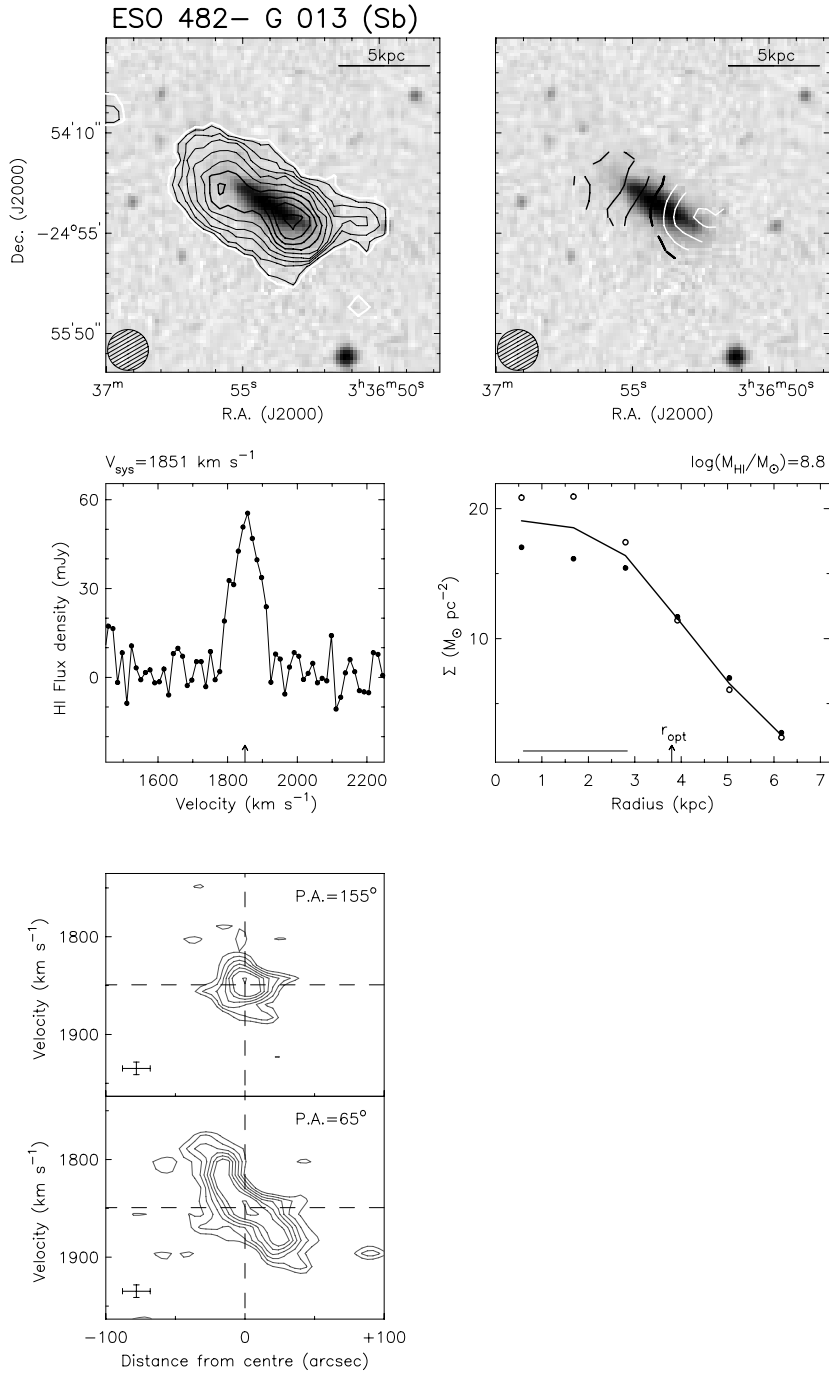
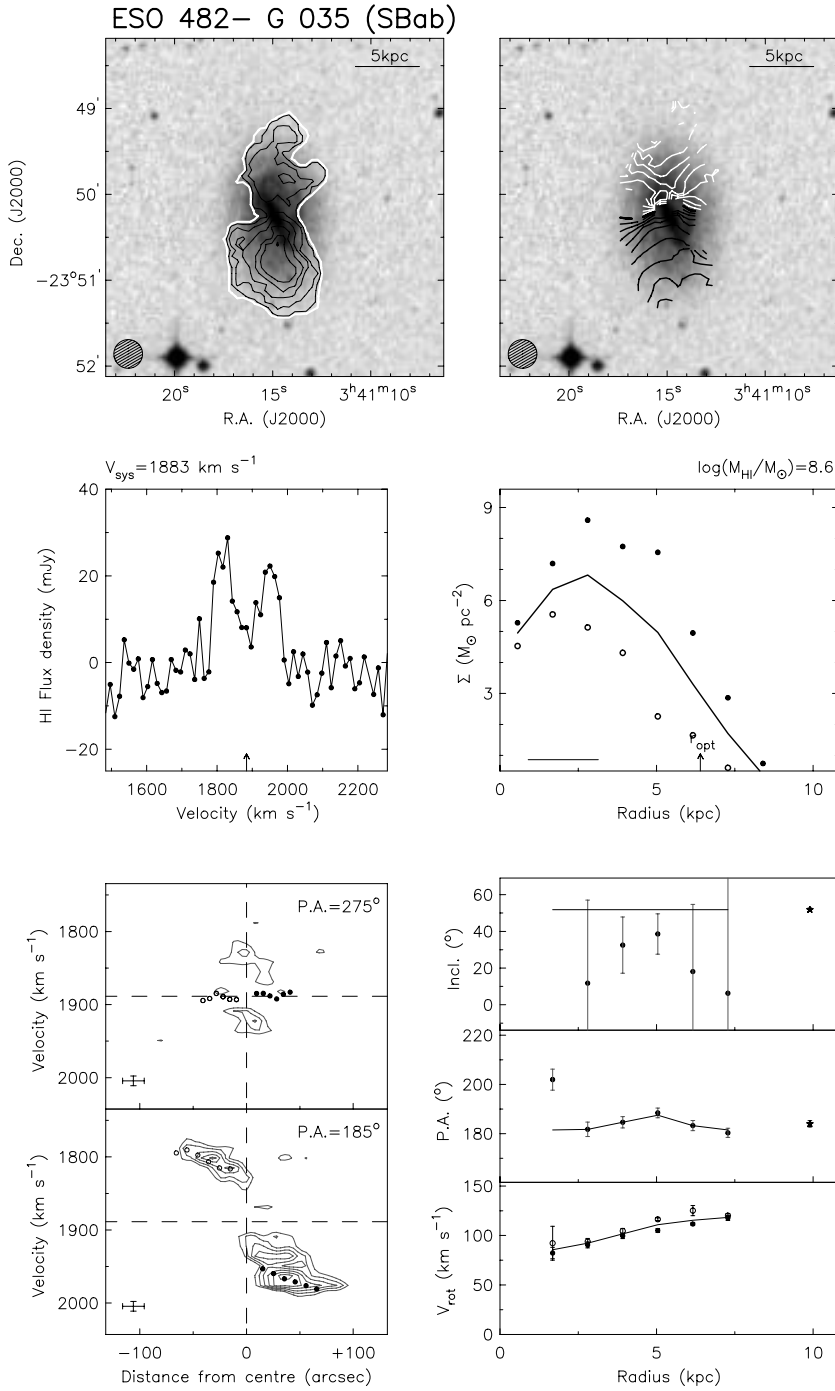
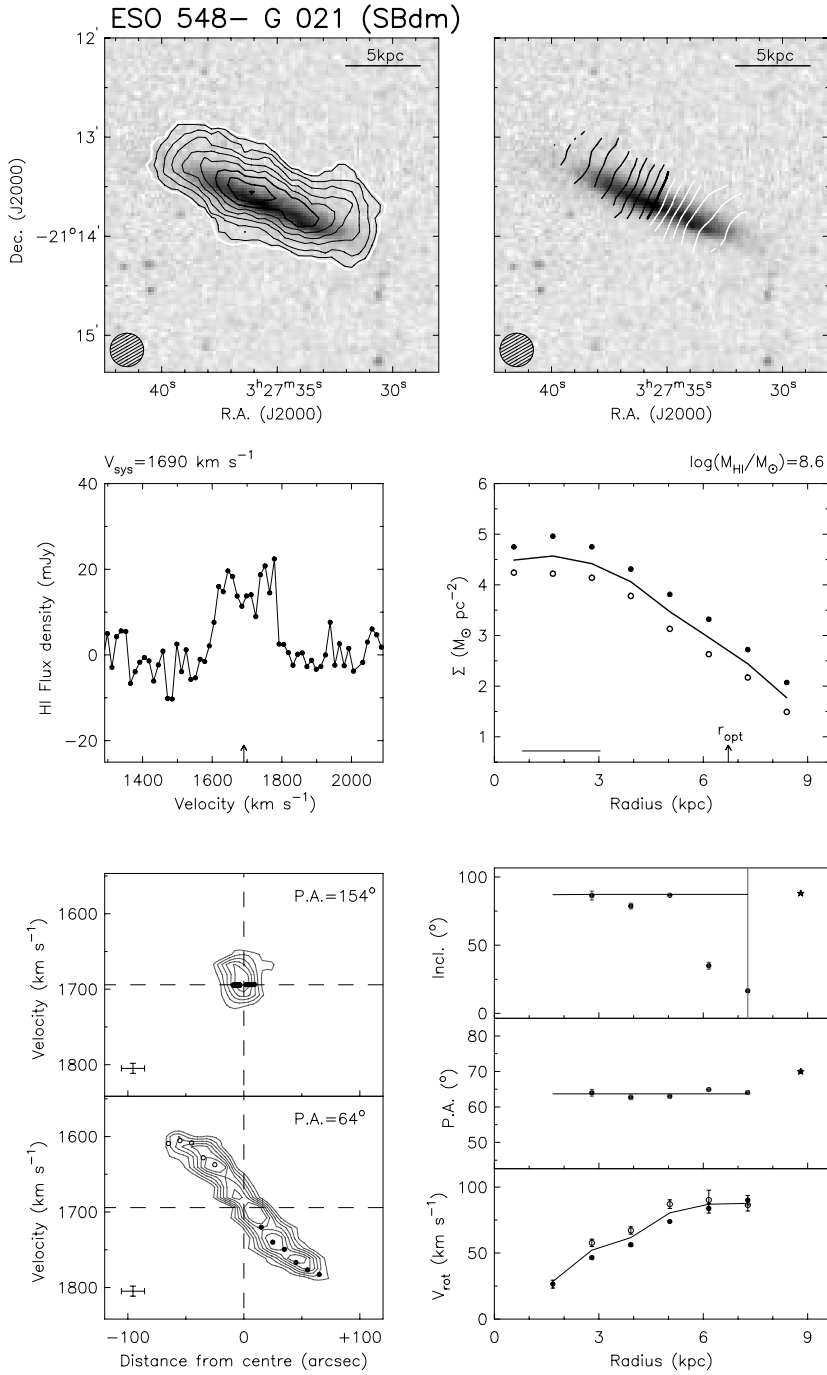
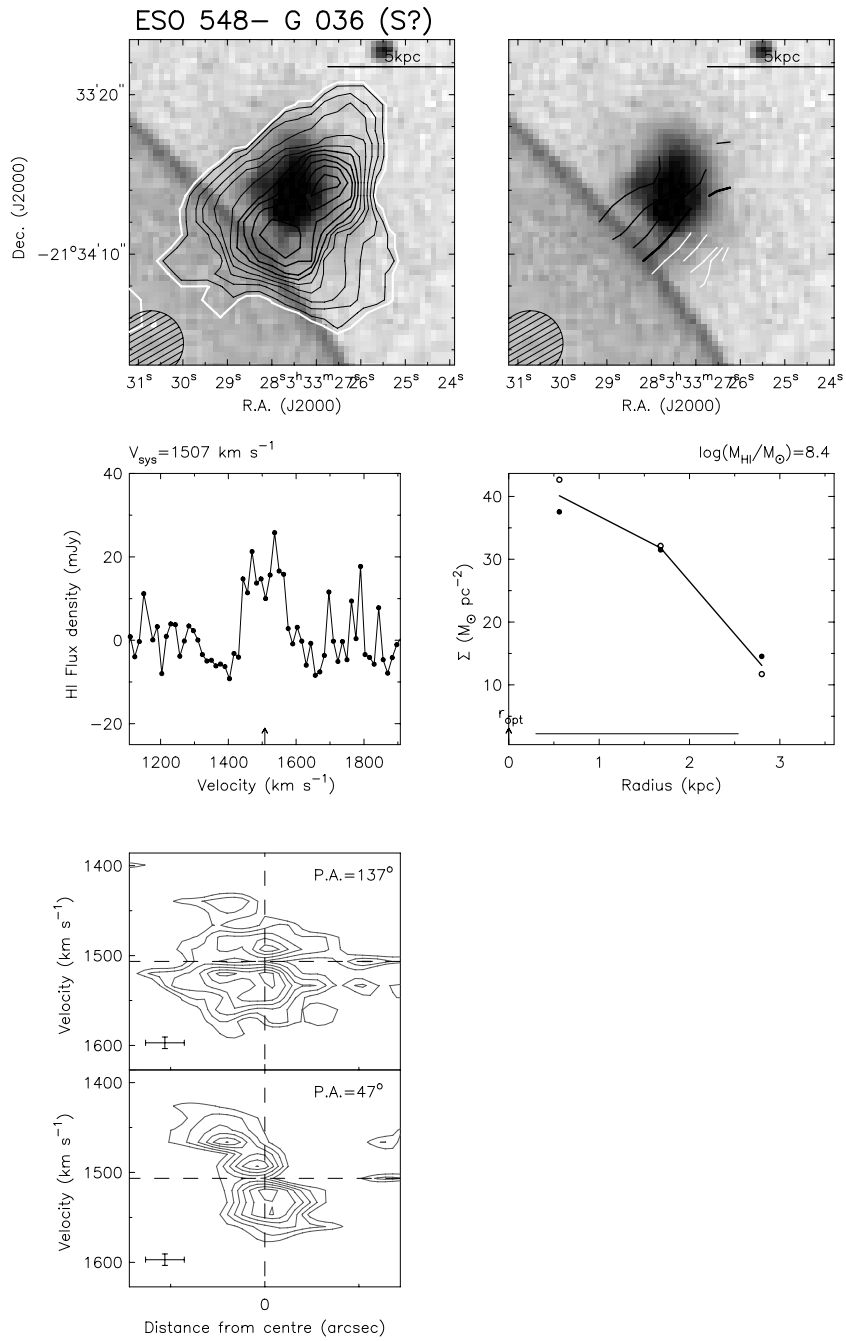


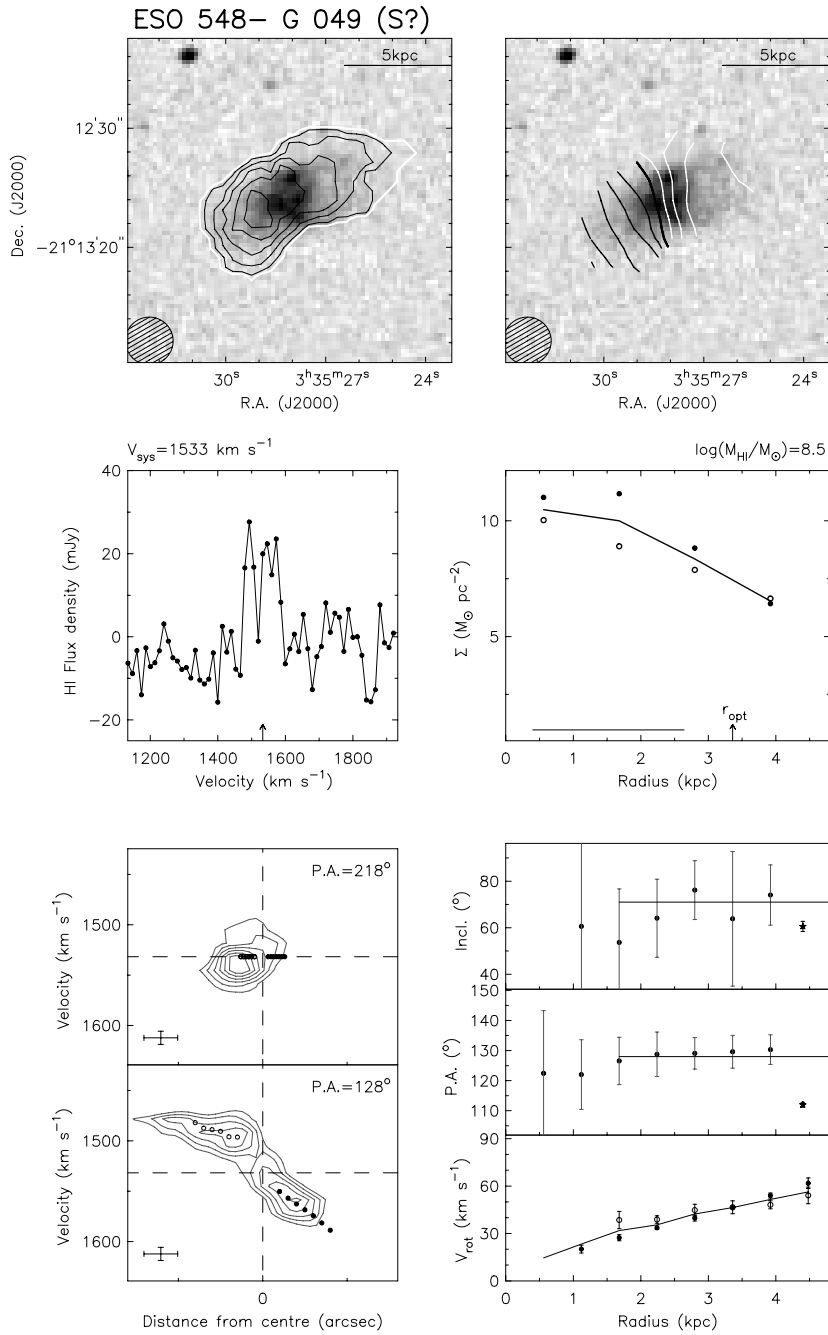
Figure 16. (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

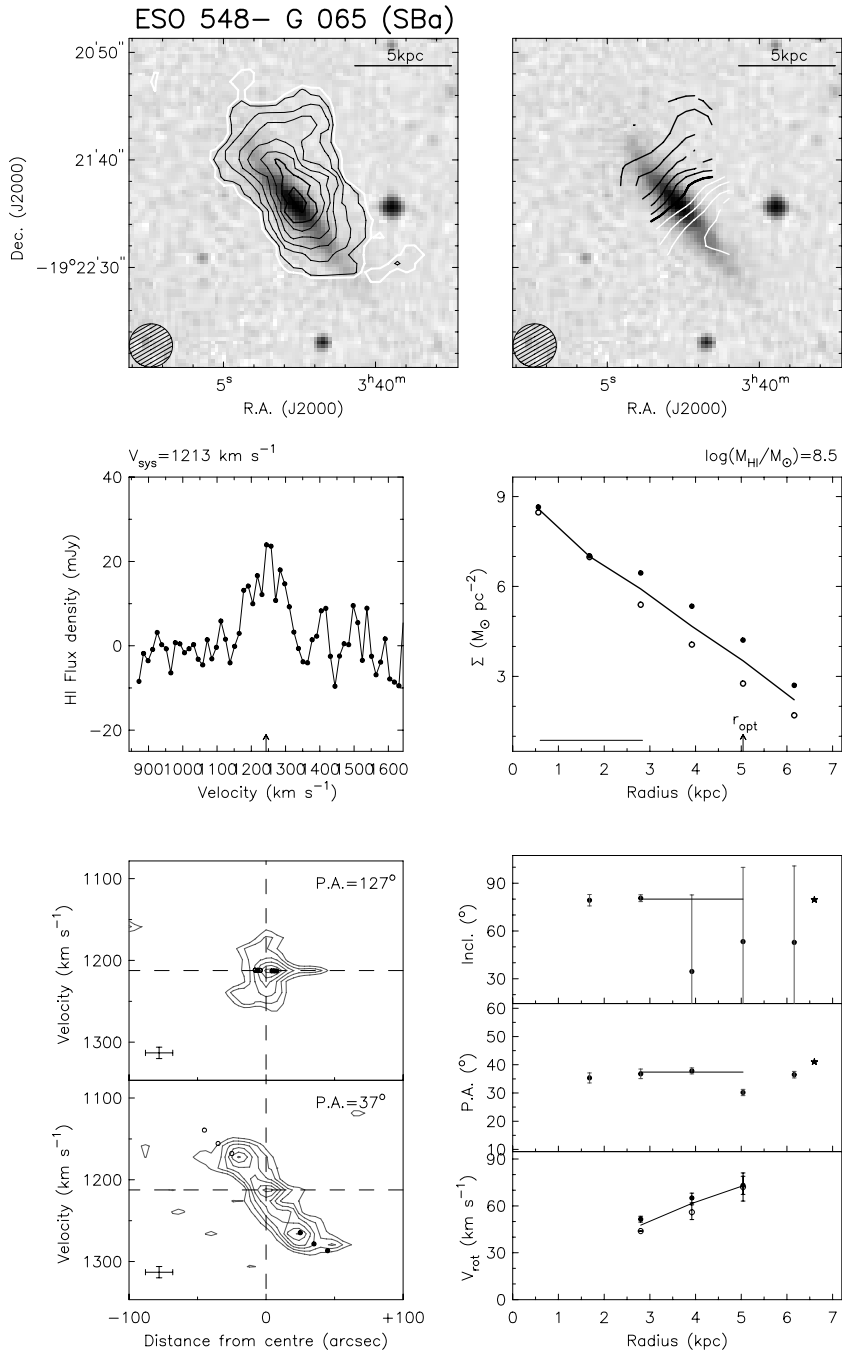
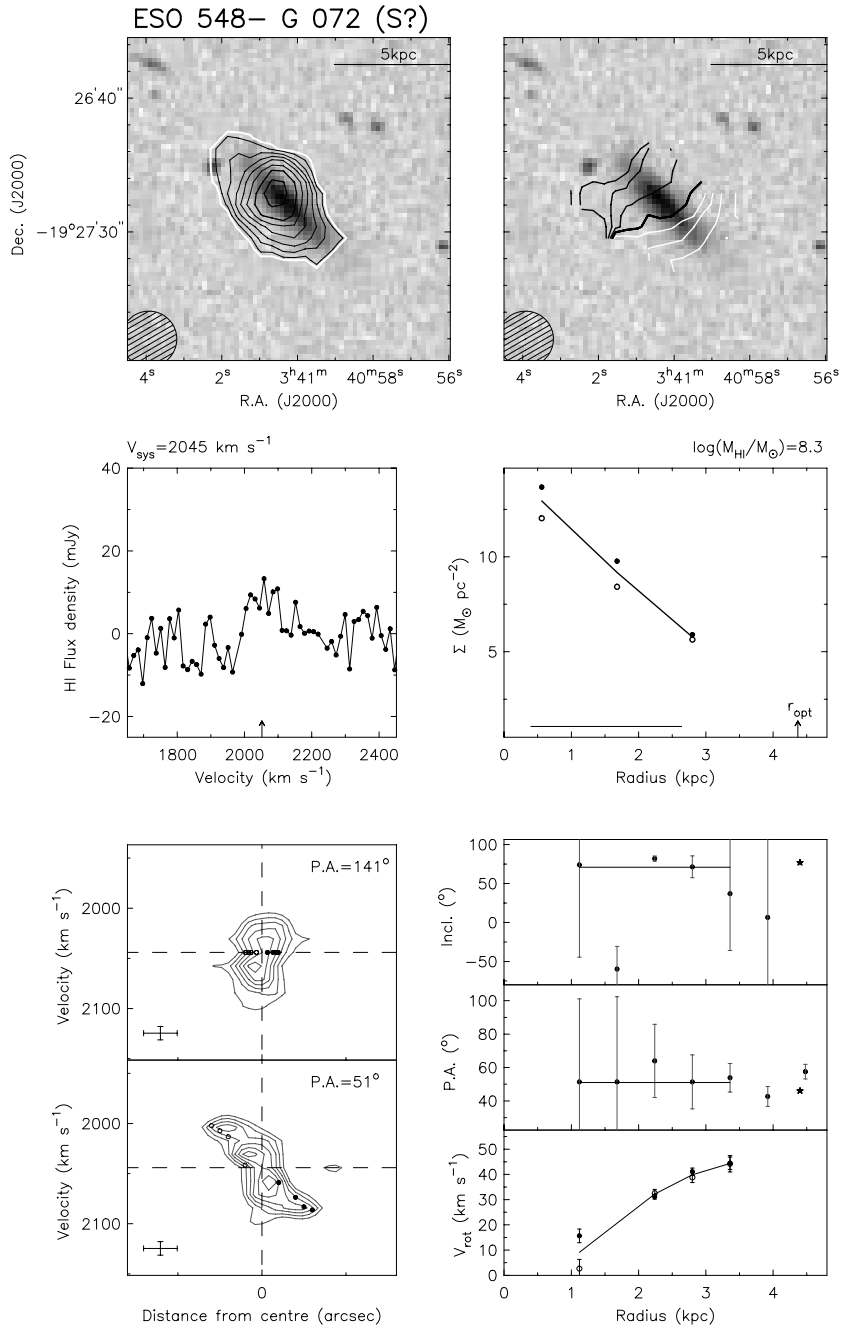
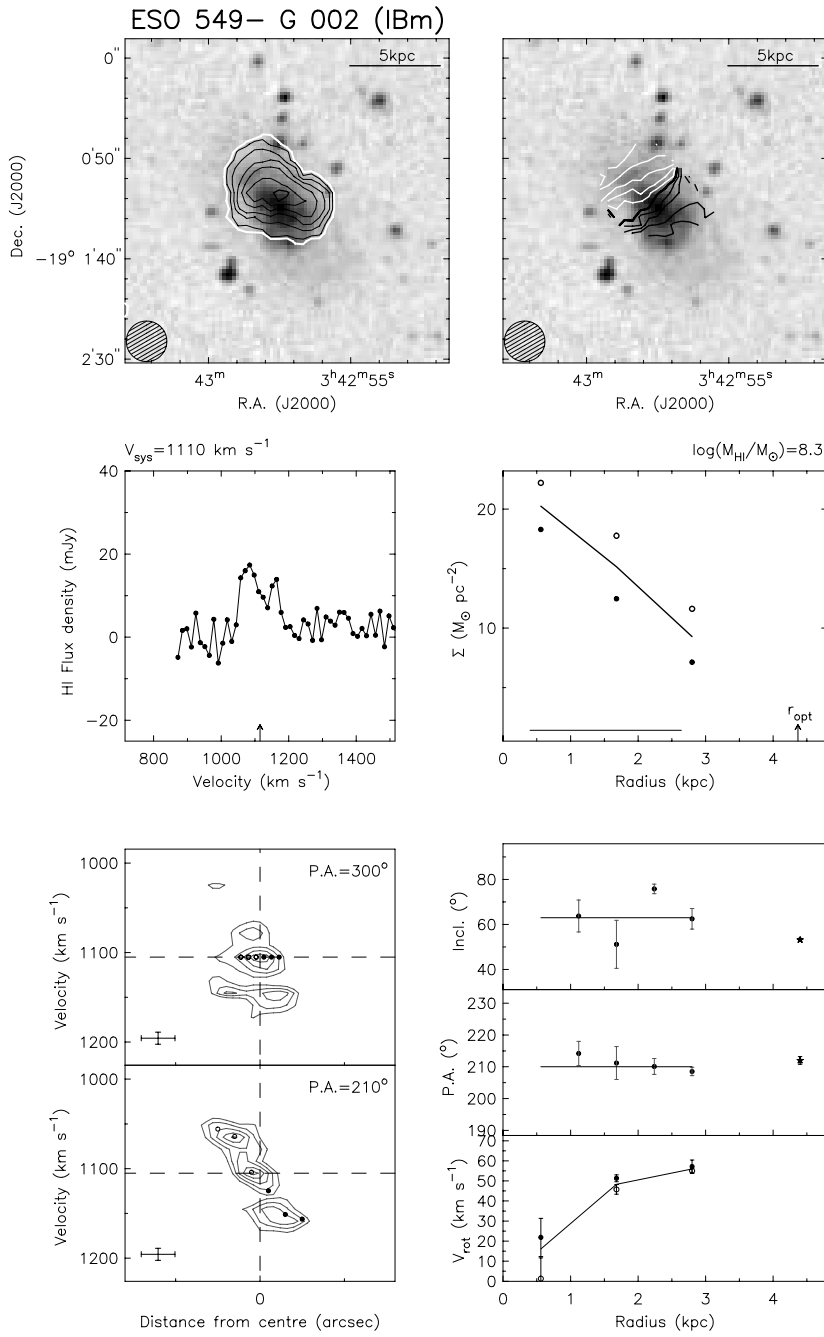
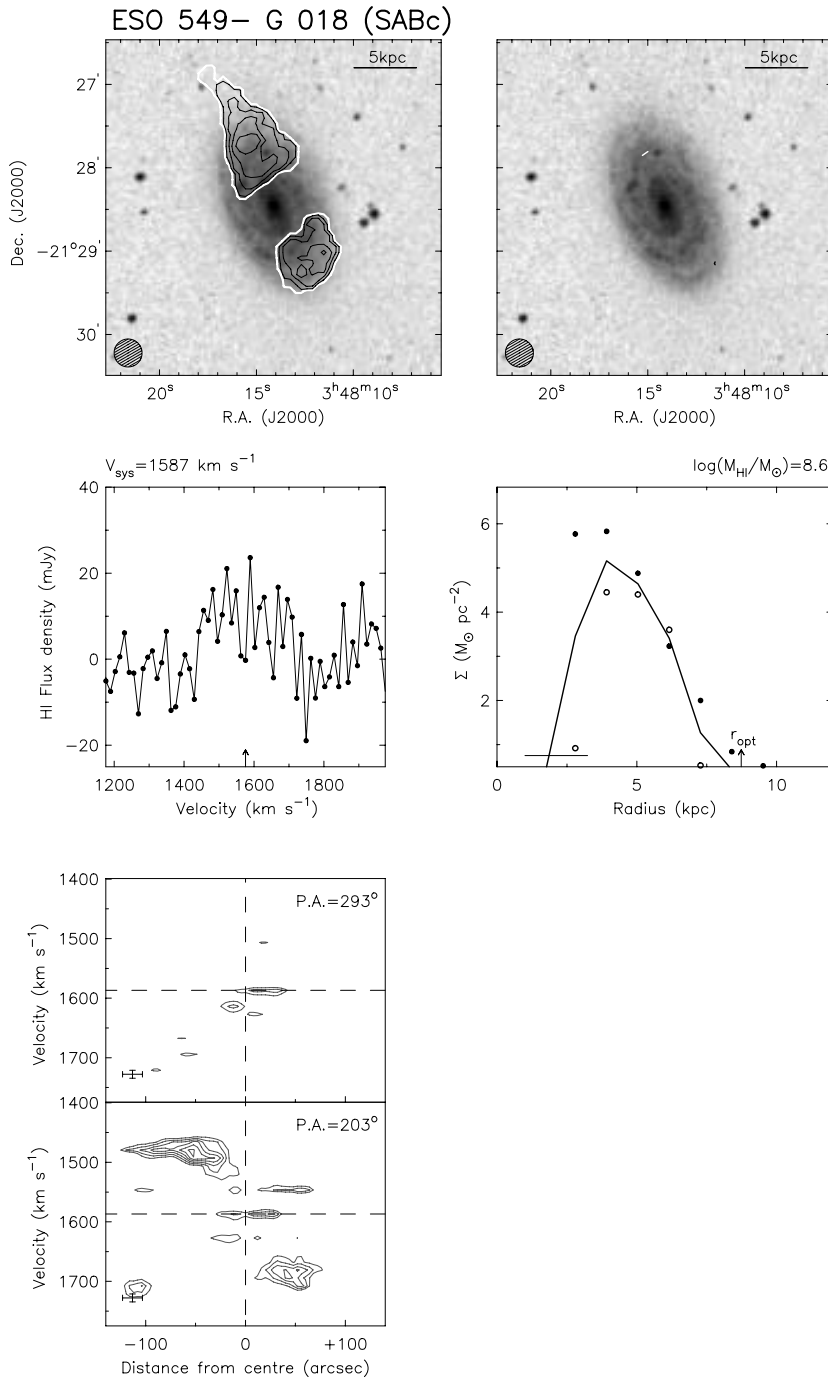
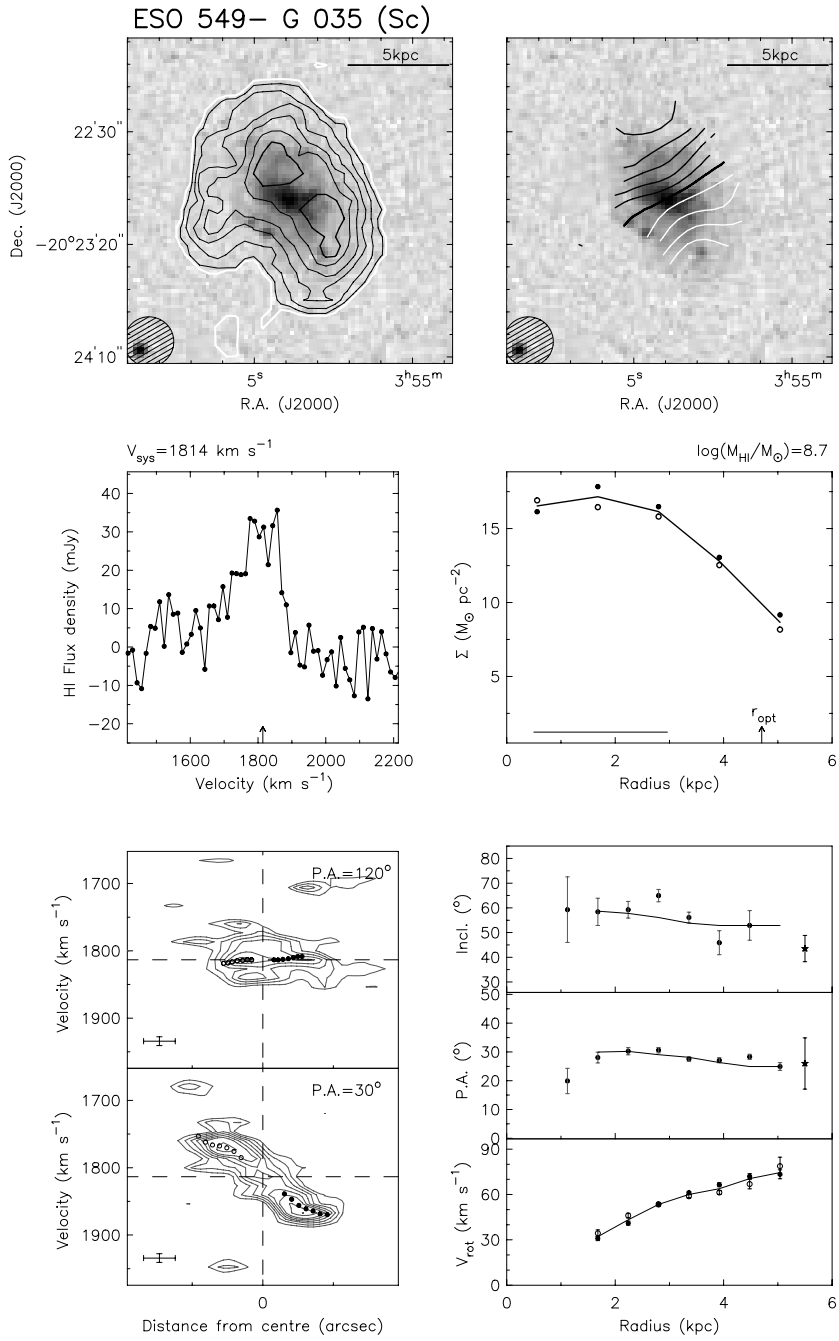


Figure 16. (Continued)

**Figure 16.** (Continued)


Figure 16. (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

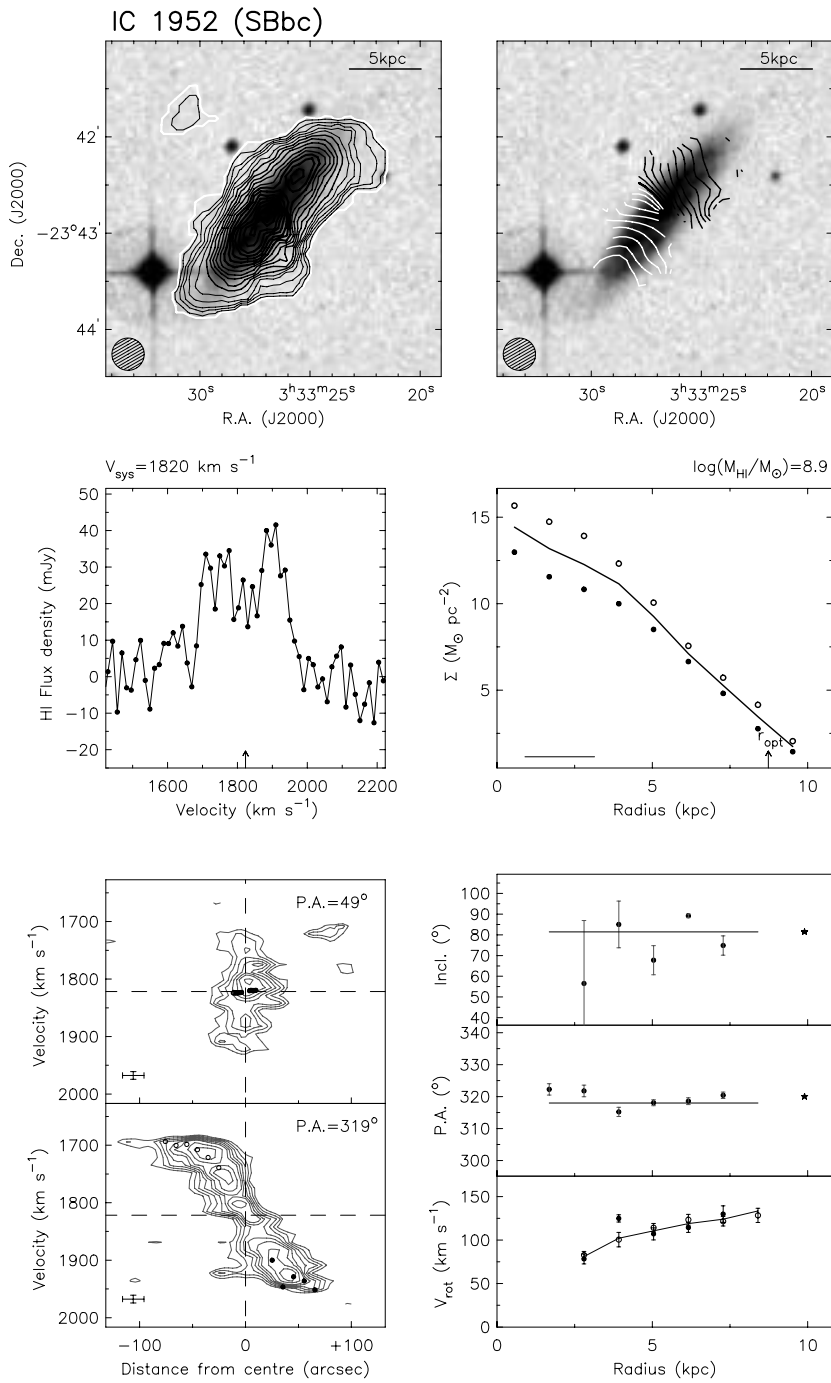
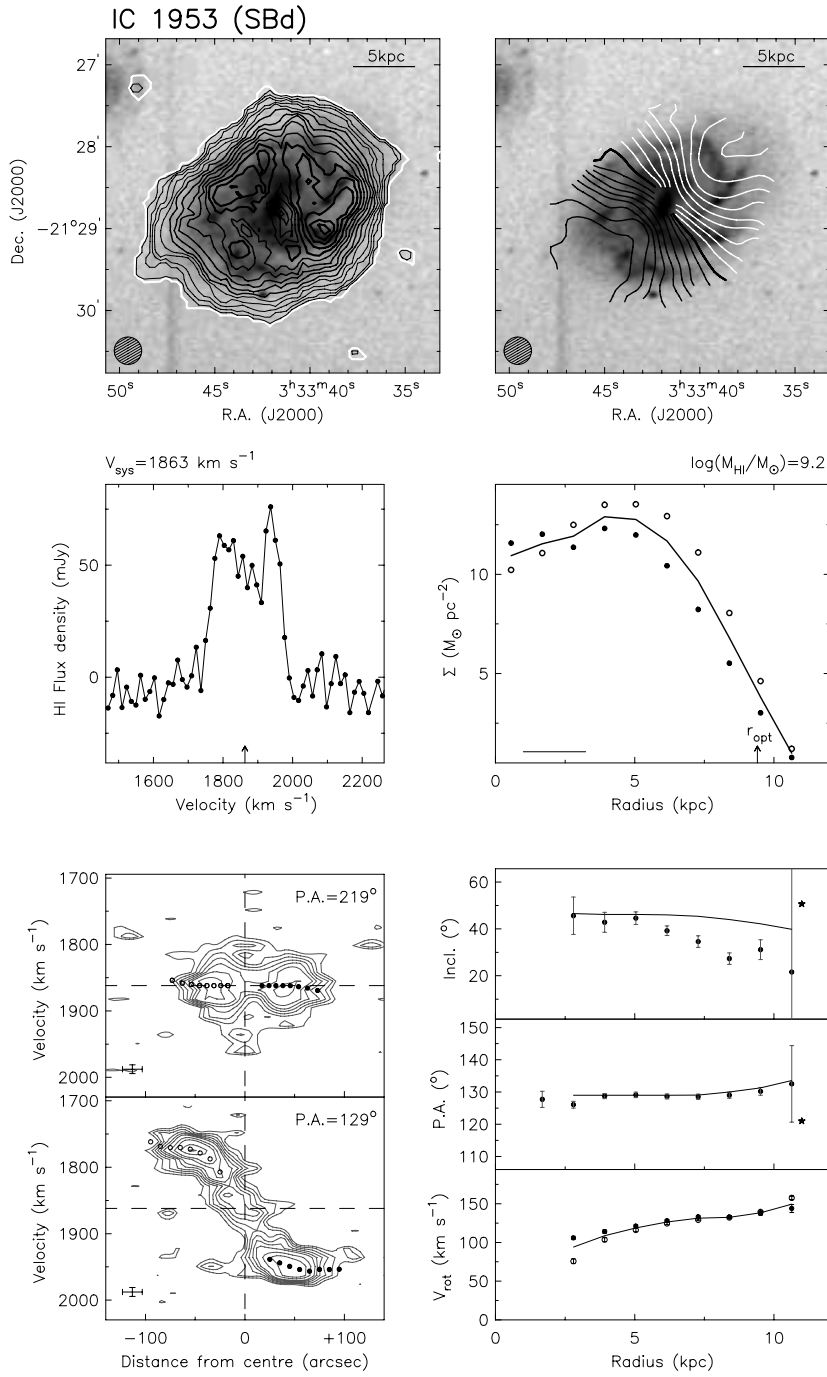


Figure 16. (Continued)

**Figure 16.** (Continued)

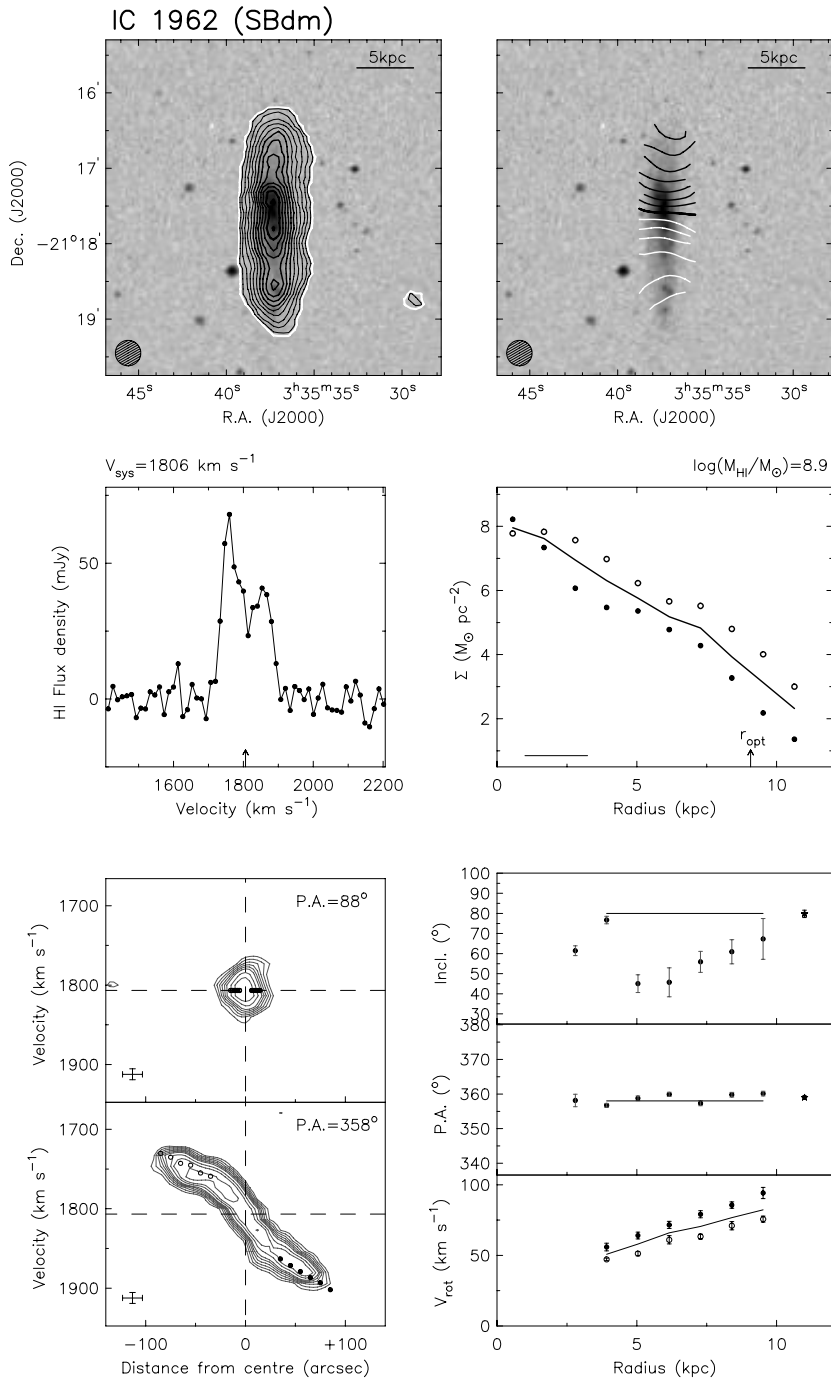
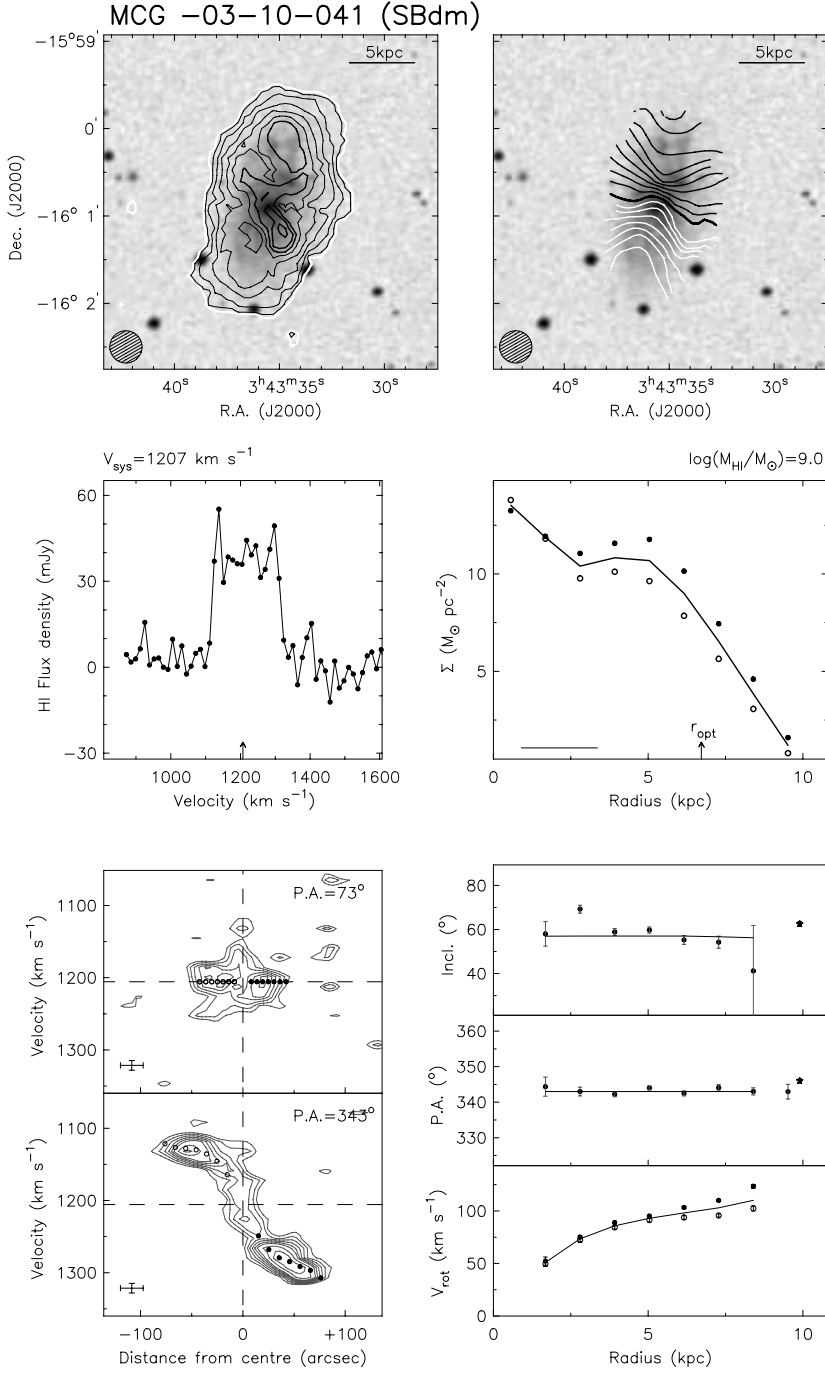
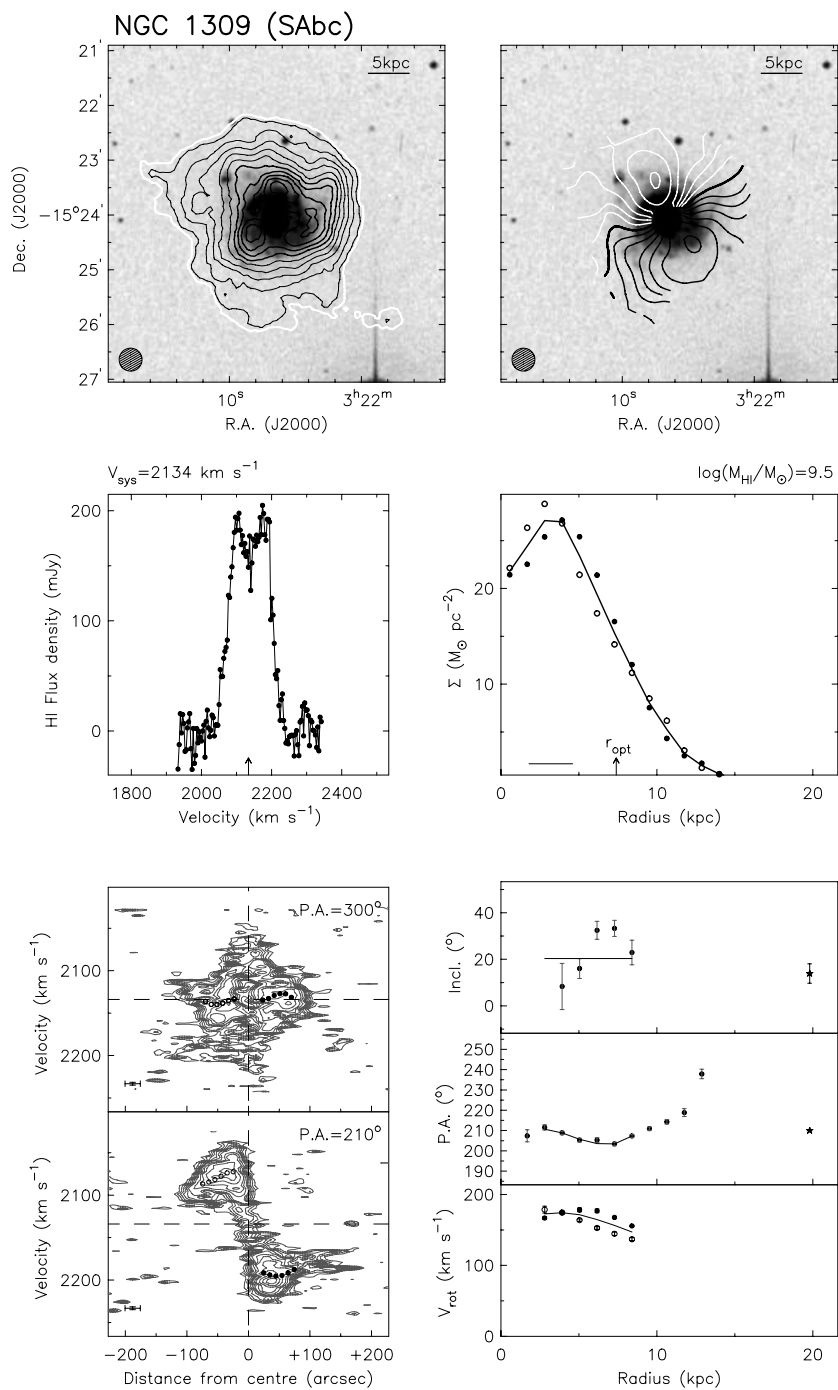
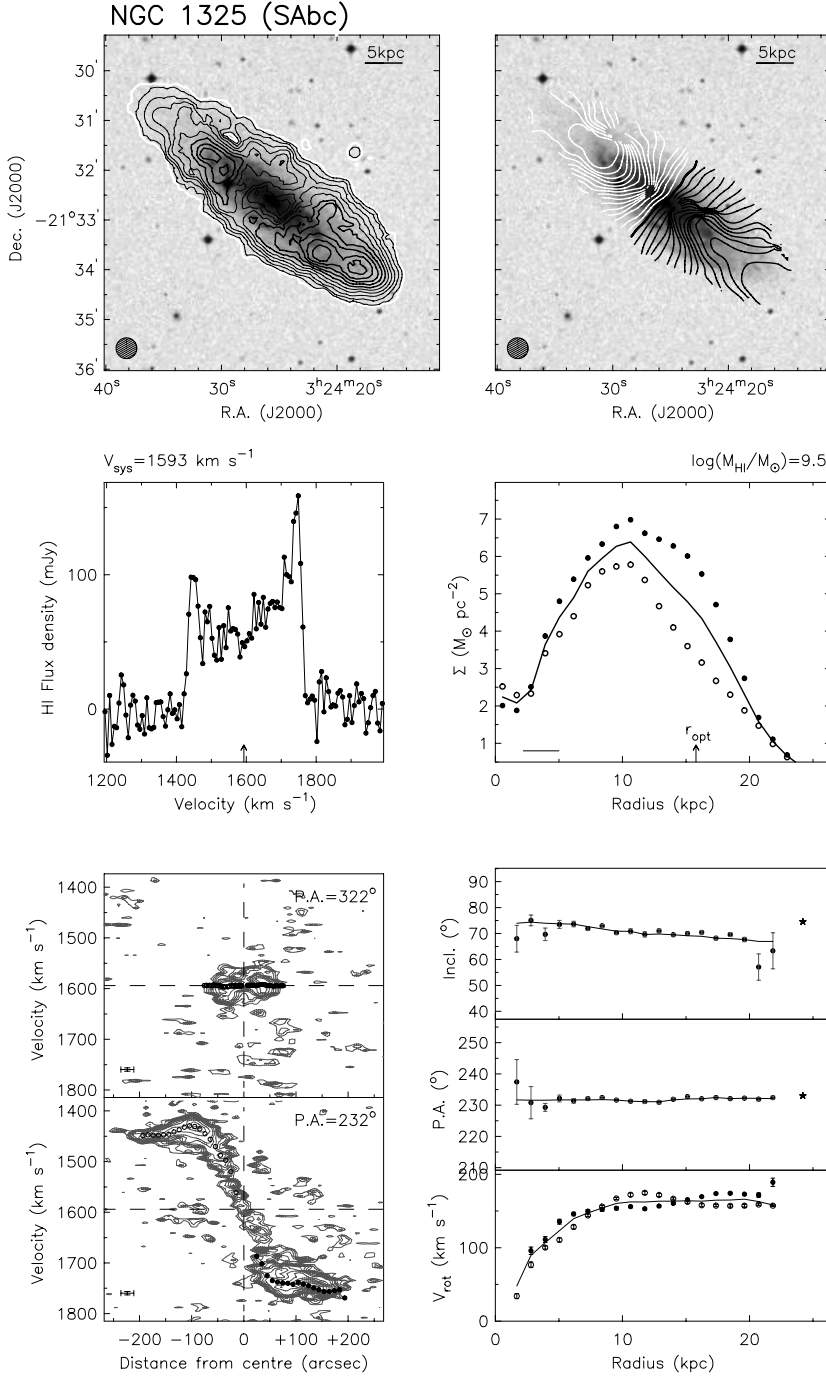
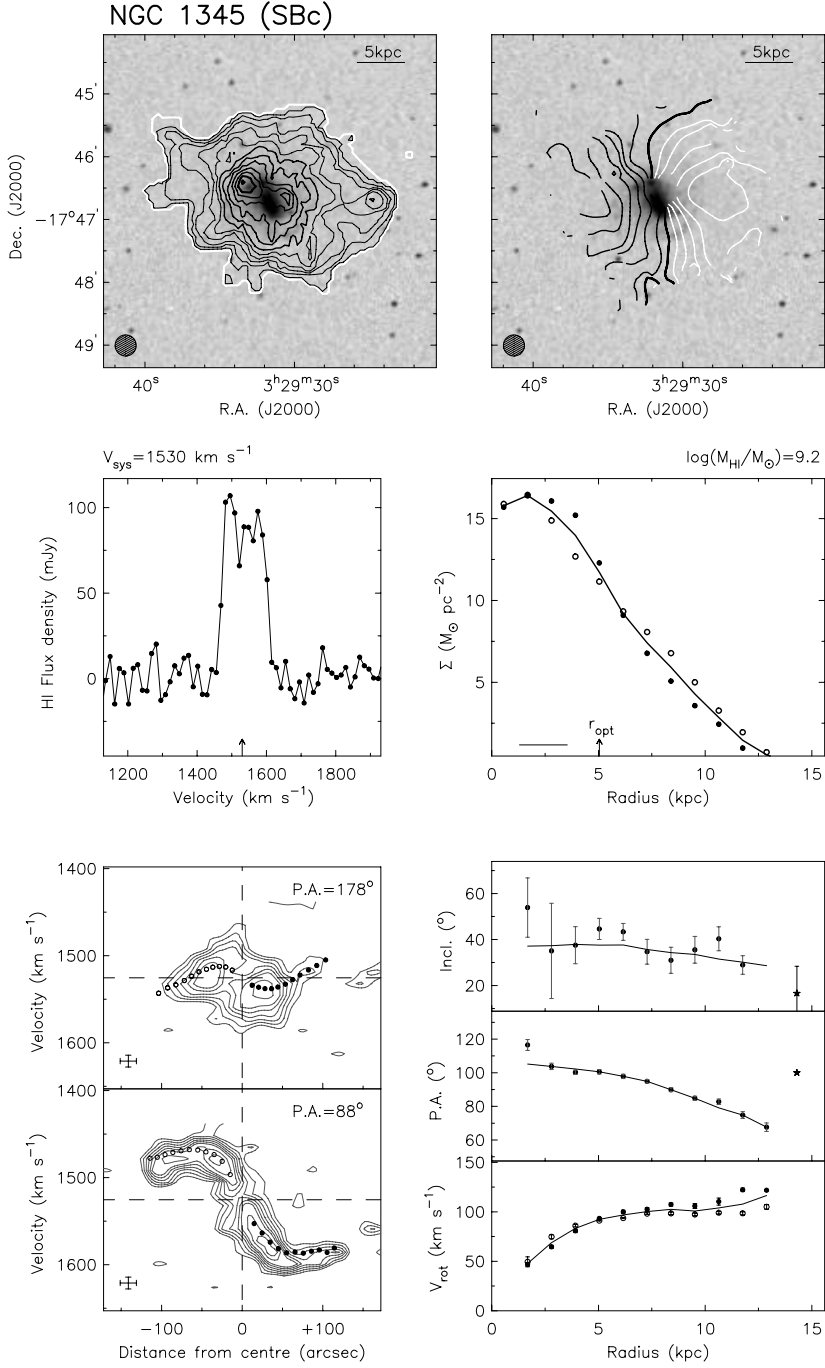


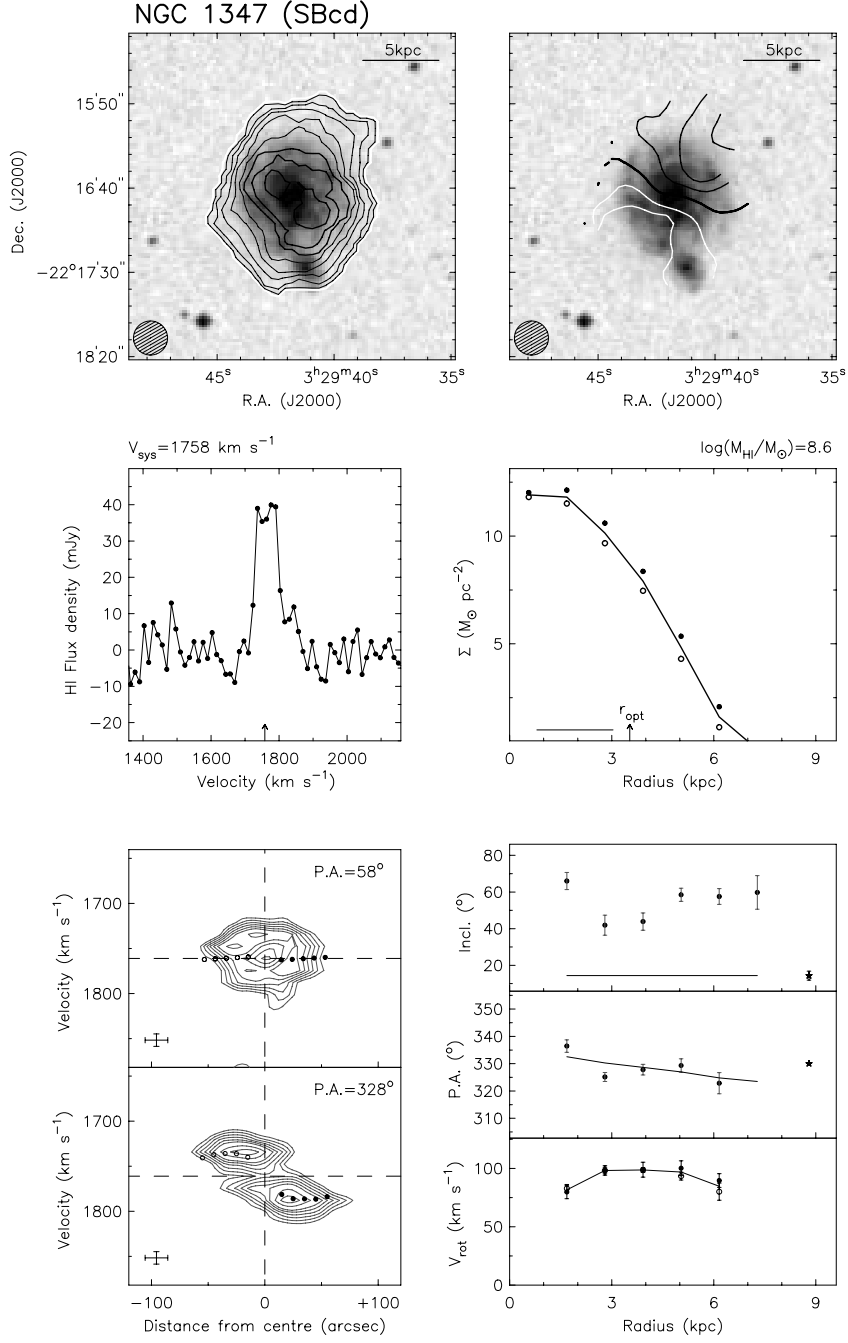
Figure 16. (Continued)

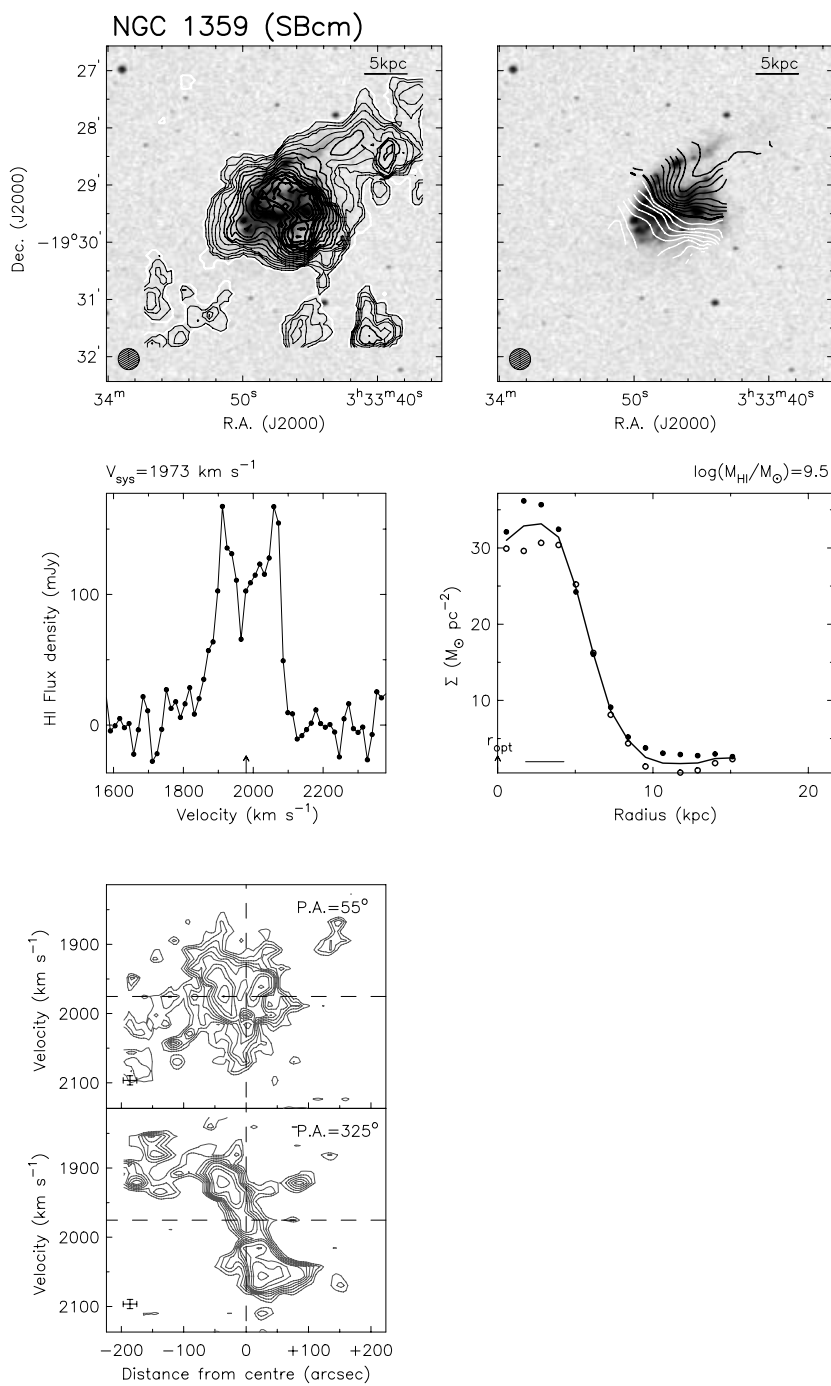
**Figure 16.** (Continued)

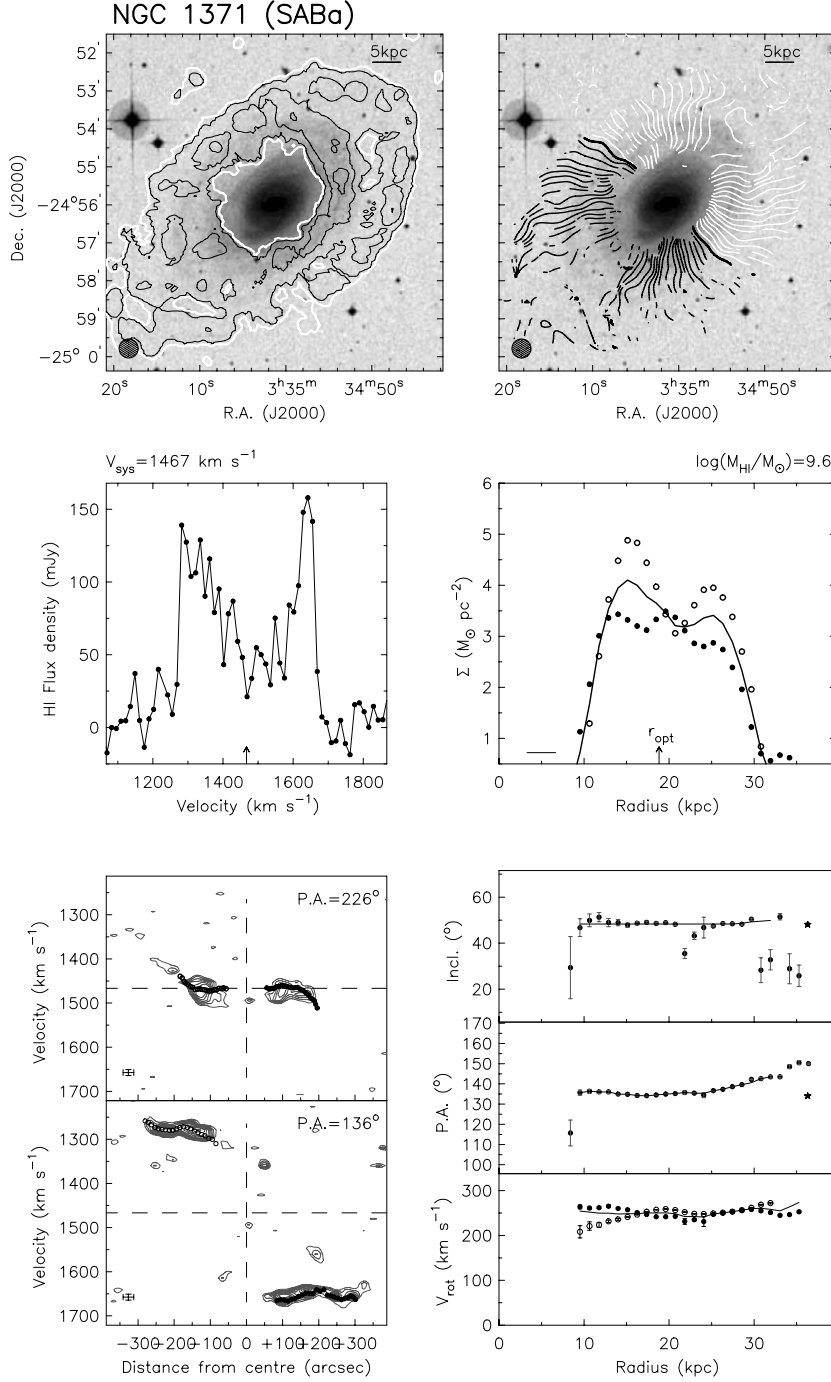
**Figure 16.** (Continued)

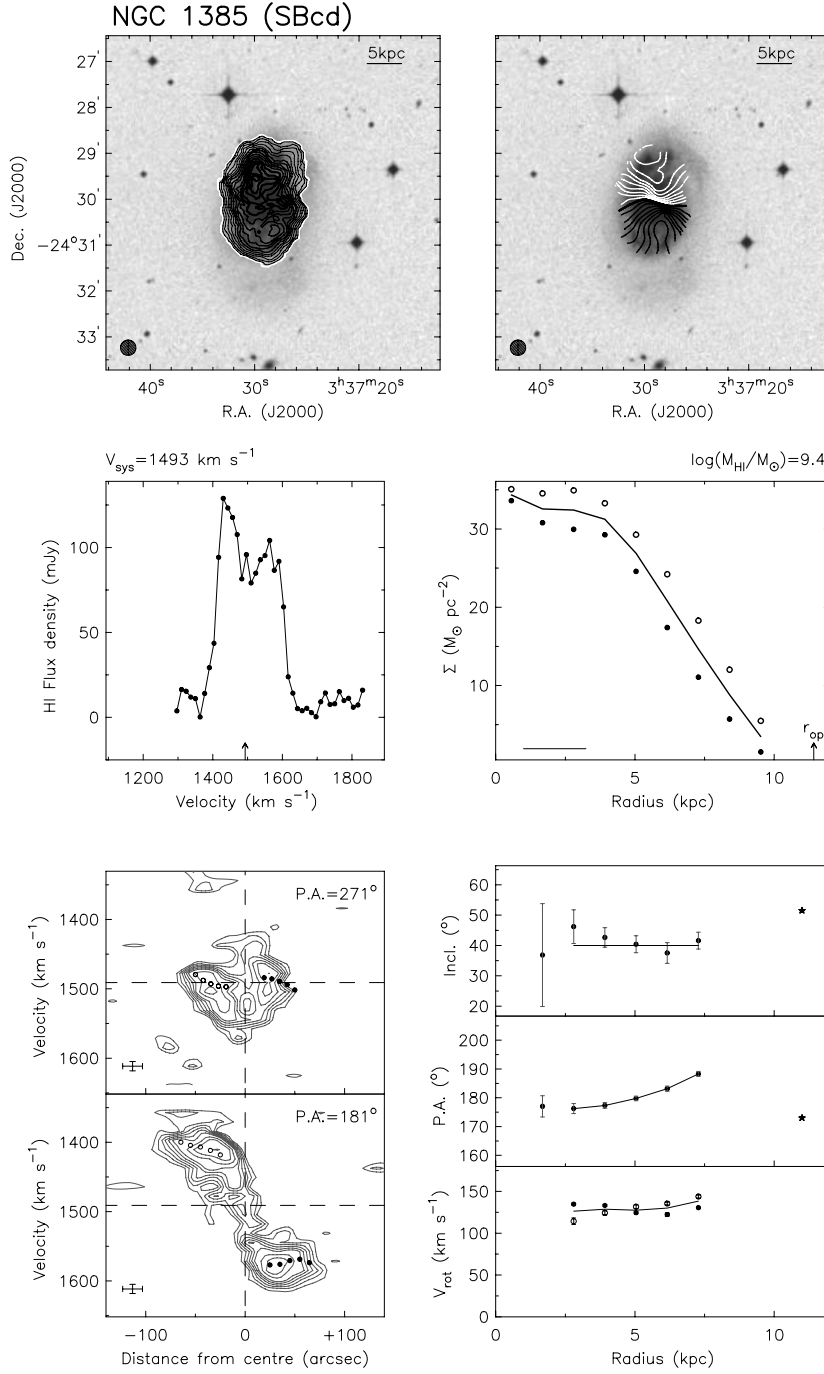
**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

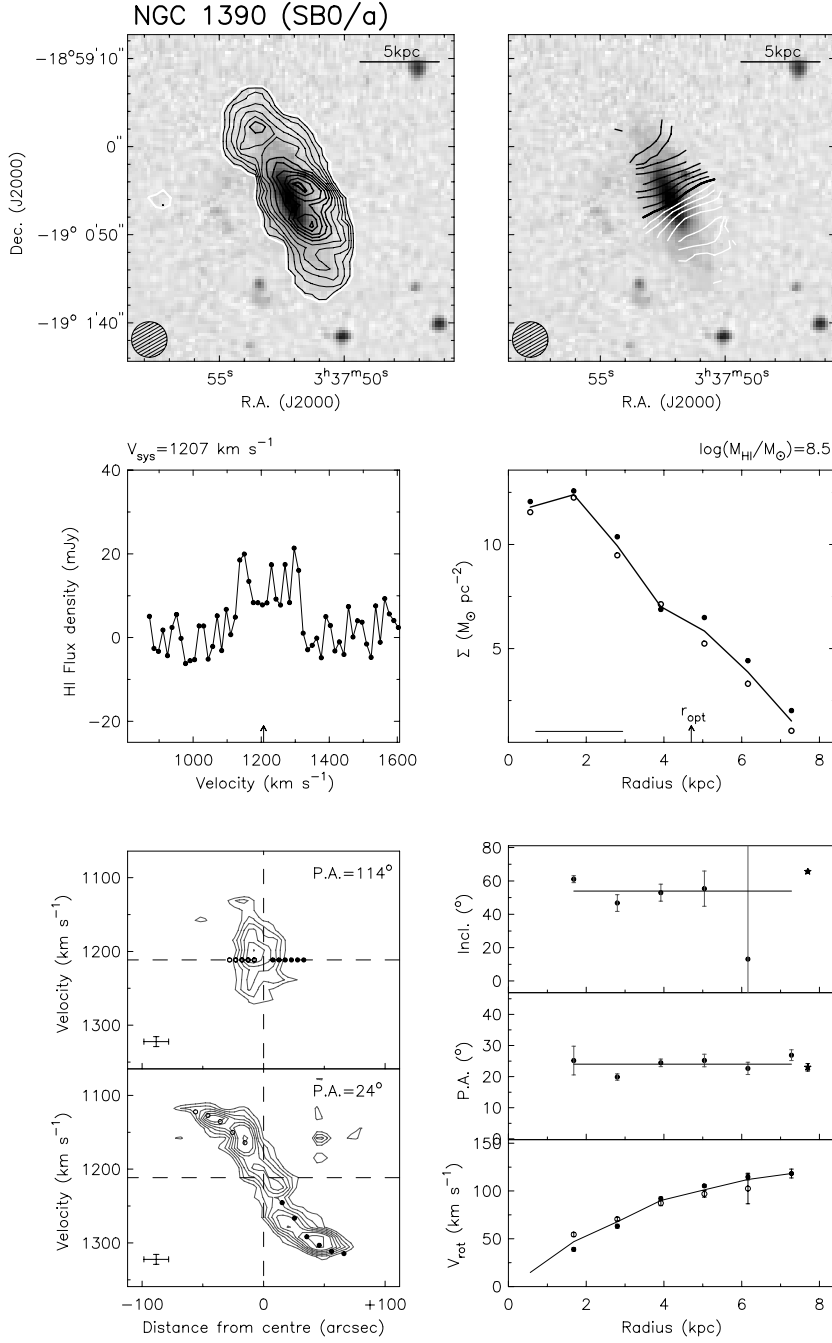
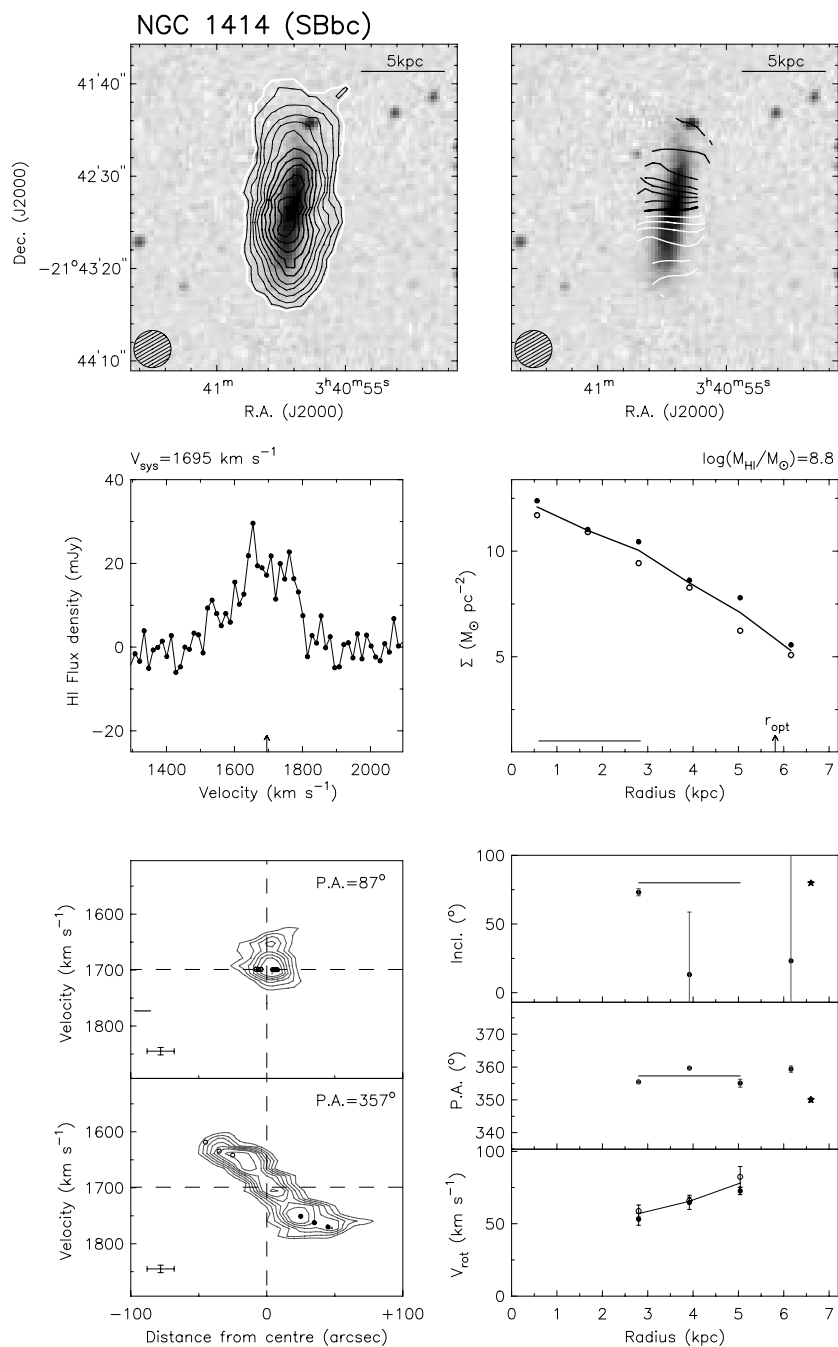
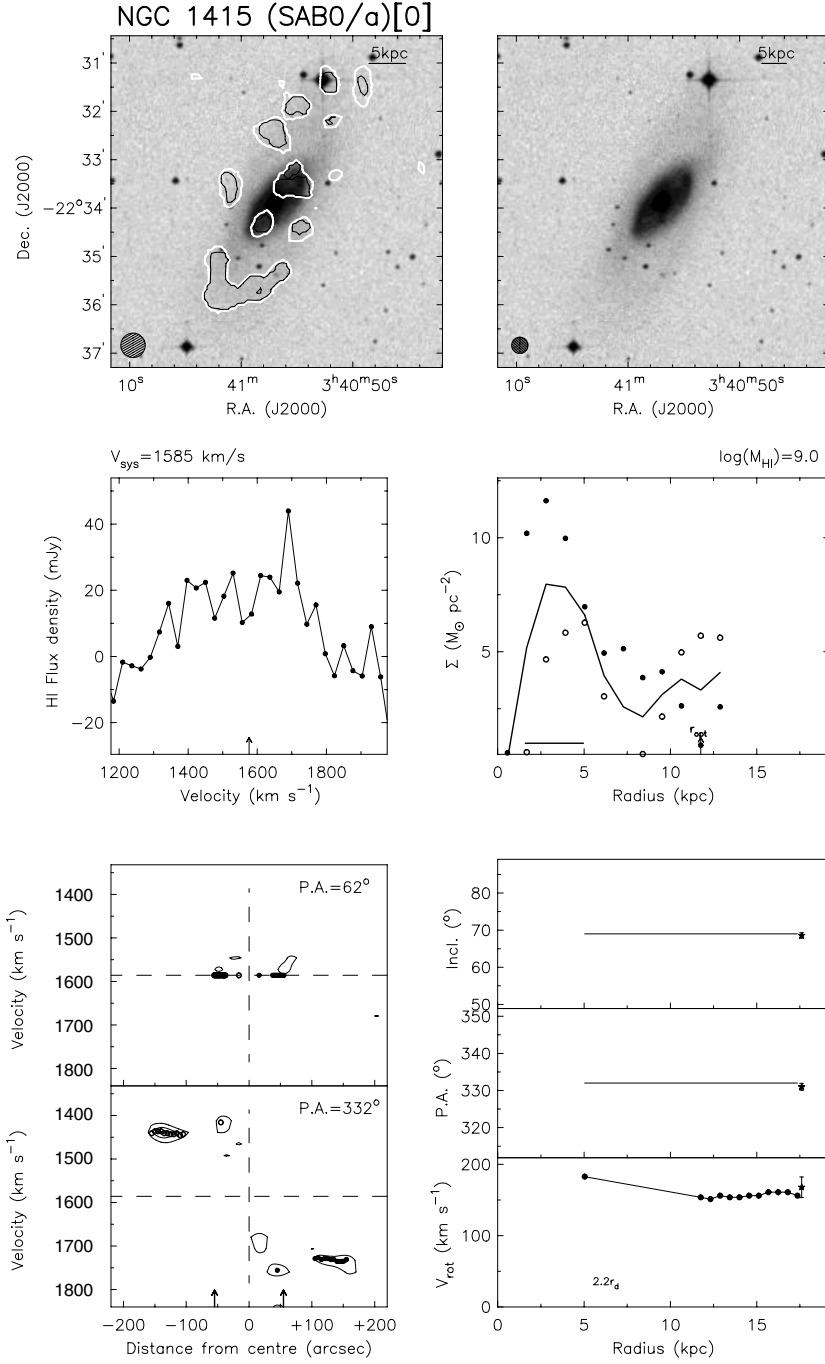
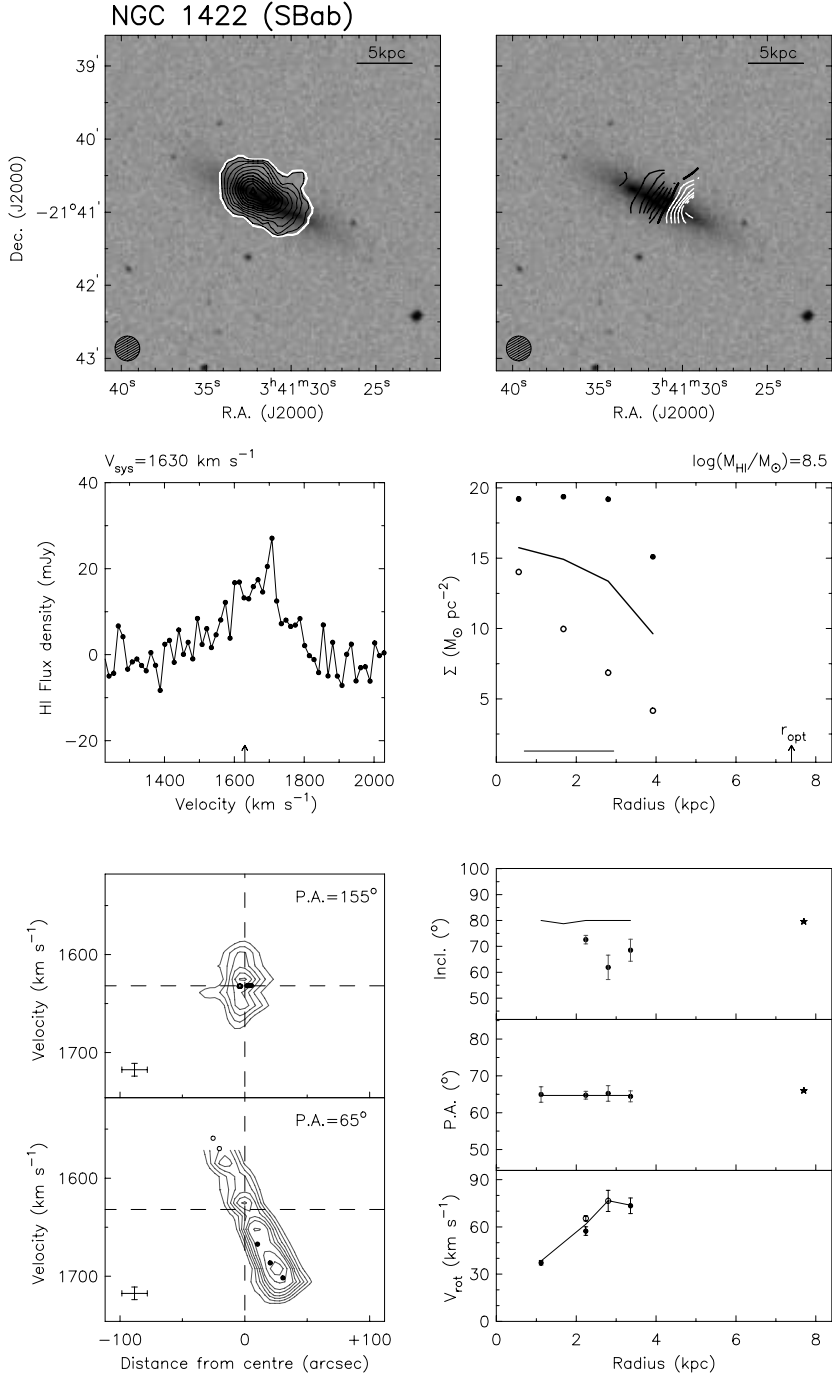


Figure 16. (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

**Figure 16.** (Continued)

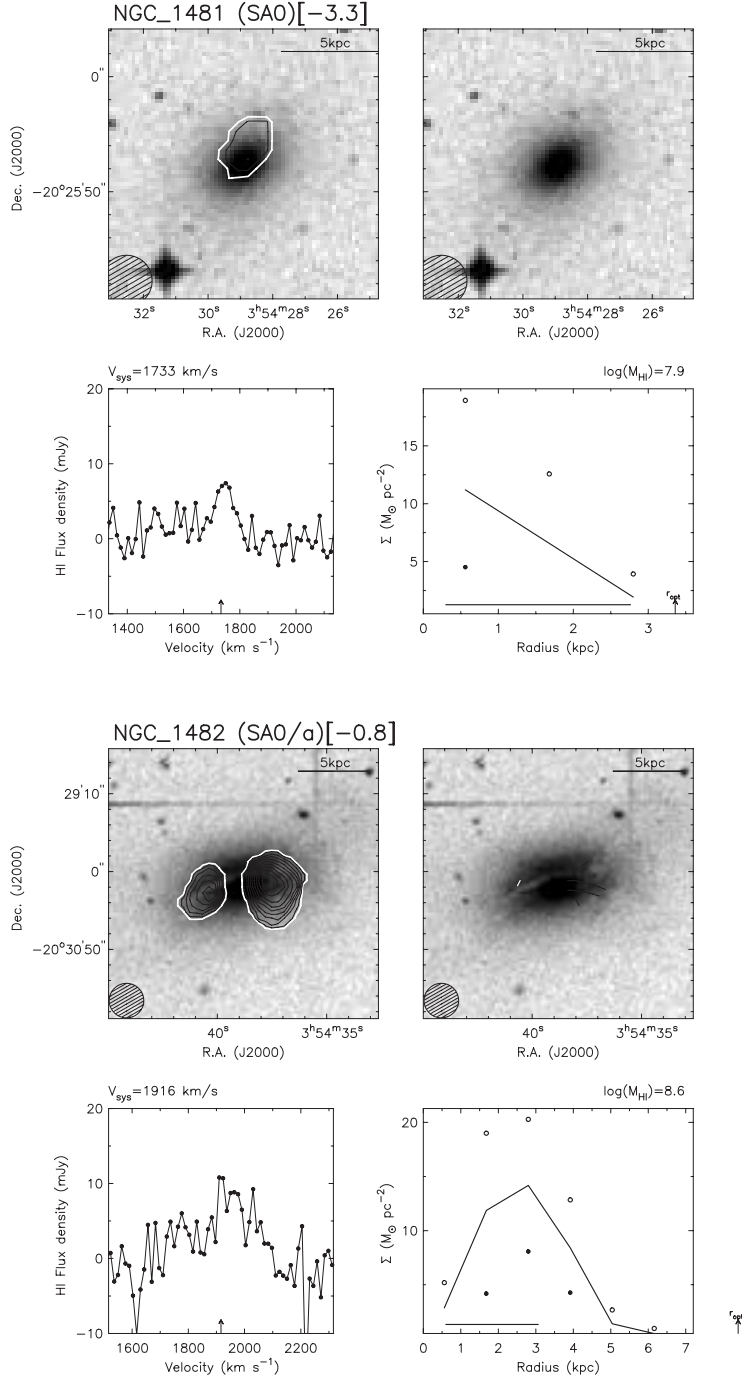
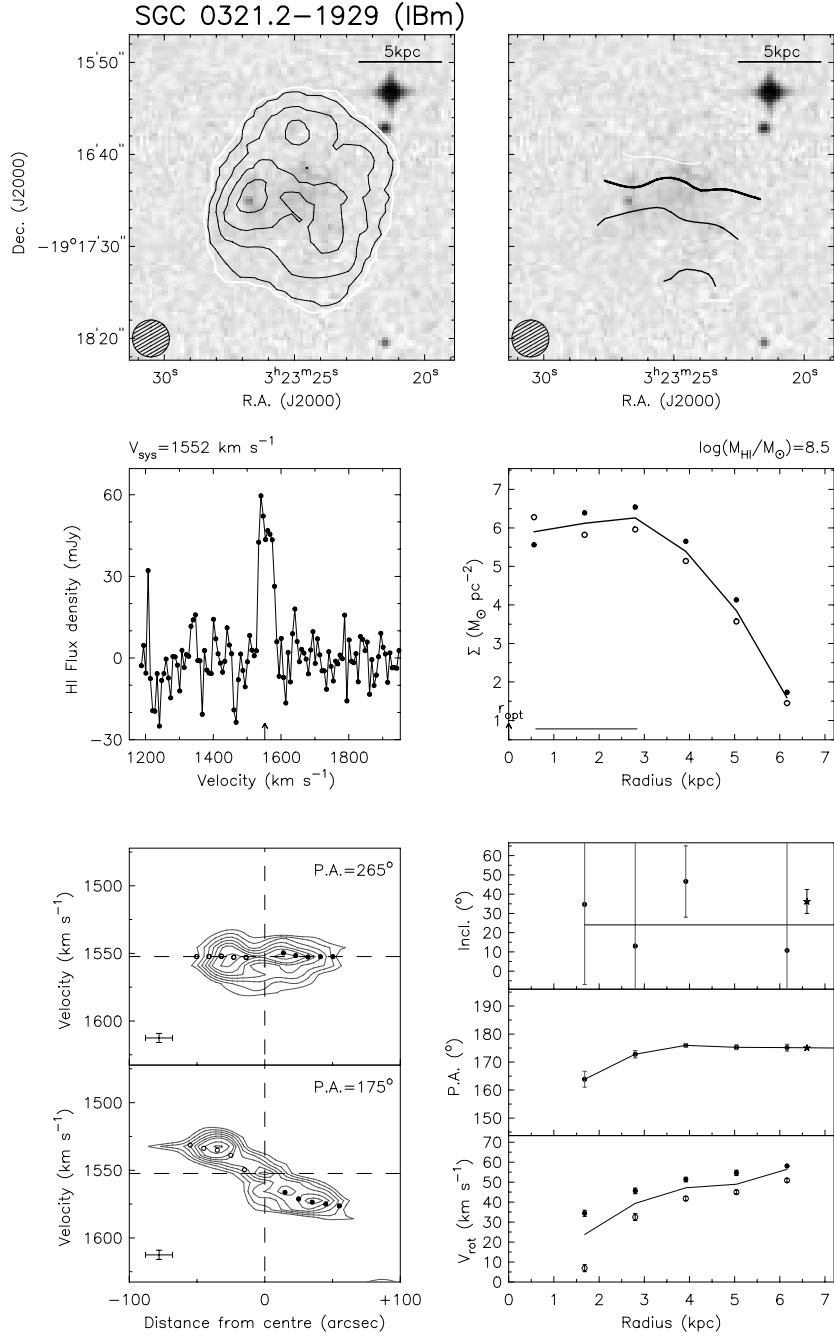


Figure 16. (Continued)

**Figure 16.** (Continued)

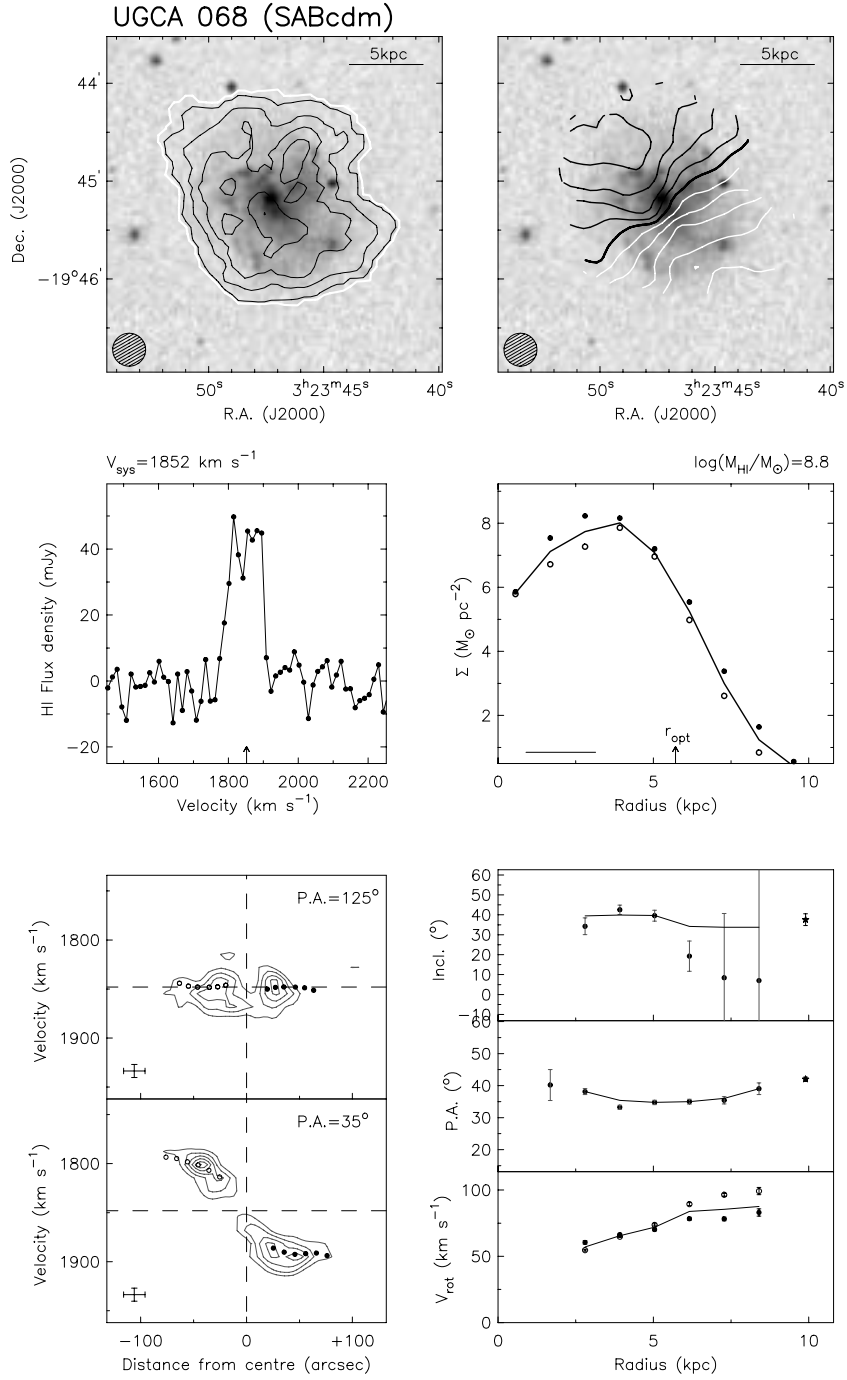
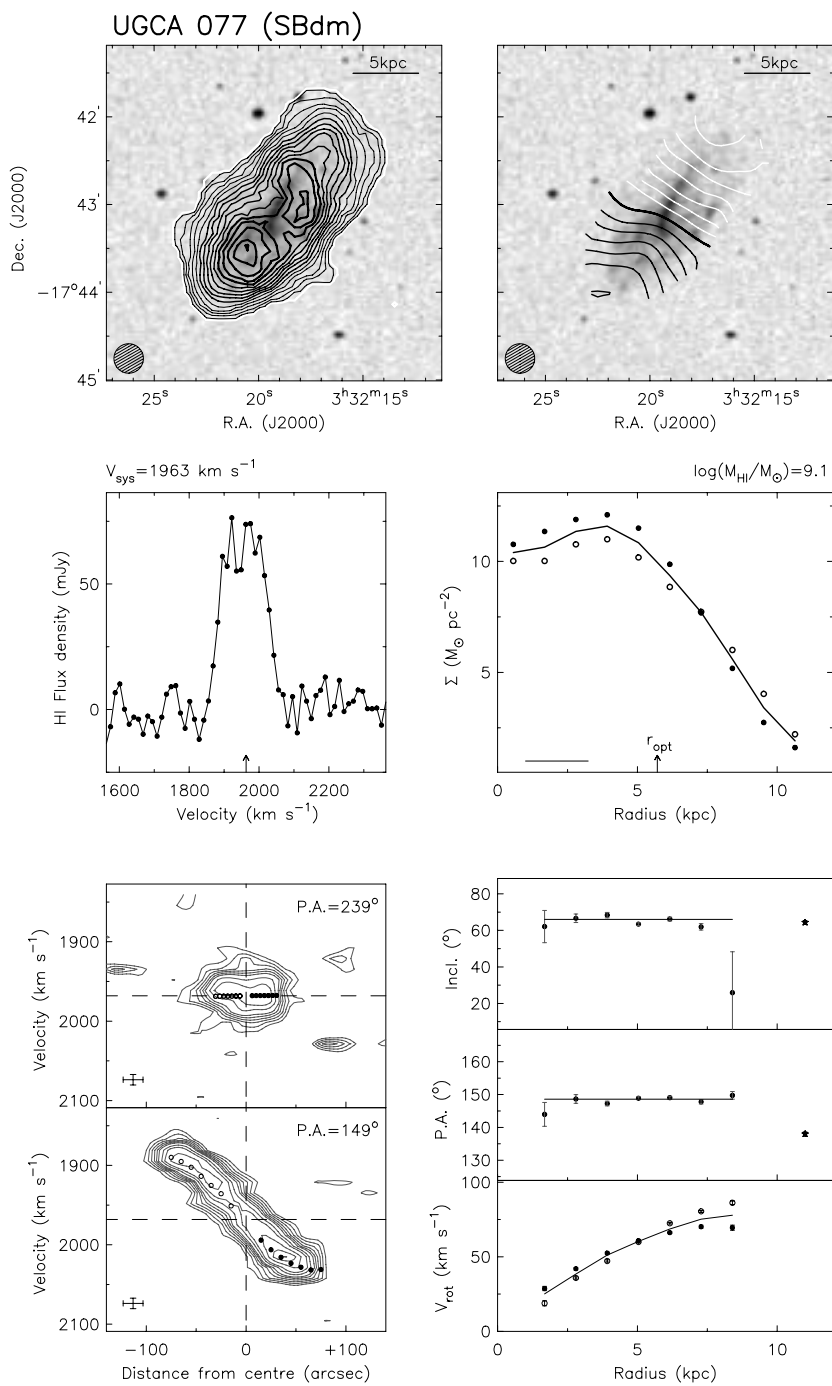


Figure 16. (Continued)

**Figure 16.** (Continued)

References

- Baker, R. H. 1933, *Ann. Harv. Coll. Obs.*, **88**, 79.
 Baker, R. H. 1936, *Ann. Harv. Coll. Obs.*, **88**, No. 6.
 Begeman, K. 1989, *Astron. Astrophys.*, **223**, 47.
 Binggeli, B., Tammann, G. A., Sandage, A. 1987, *Astron. J.*, **94**, 251.
 Bottinelli, L., Gouguenheim, L., Fouque, P., Paturel, G. 1990, *Astron. Astrophys. Suppl.*, **82**, 391.
 Bravo-Alfaro, H., Cayatte, V., van Gorkom, J. H., Balkowski, C. 2000, *Astron. J.*, **119**, 580.
 Broeils, A. H., Rhee, M.-H. 1997, *Astron. Astrophys.*, **324**, 877.
 Cayatte, V., van Gorkom, J. H., Balkowski, C., Kotanyi, C. 1990, *Astron. J.*, **100**, 604.
 Cowie, L. L., Songaila, A. 1977, *Nature*, **266**, 501.
 Curtis, H. D. 1918, *Publ. Lick Obs.*, **13**, 9.
 Davies, R. D., Lewis, B. M. 1973, *Mon. Not. R. Astron. Soc.*, **165**, 231.
 da Costa, L. N., Pellegrini, P. S., Sargent, W. L. *et al.* 1988, *Astrophys. J.*, **327**, 544.
 de Vaucouleurs, G. In *Galaxies and the Universe*, (eds) Sandage, A., Sandage, M., Kristian, J., Univ. of Chicago, Chicago, p. 557.
 Dressler, A. 1980, *Astrophys. J.*, **236**, 351.
 Dressler, A., Oemler, A. Jr., Couch, W. J. *et al.* 1997, *Astrophys. J.*, **490**, 577.
 Fasano, G., Poggianti, B. M., Couch, W. J. *et al.* 2000, *Astrophys. J.*, **542**, 673.
 Federspiel, M., Tammann, G. A., Sandage, A. 1998, *Astrophys. J.*, **495**, 115.
 Ferguson, H. C. 1989, *Astroph. Space Sci.*, **157**, 227.
 Giovanelli, R., Haynes, M. P. 1985, *Astrophys. J.*, **292**, 404.
 Gooch, R., 1996 In *Astronomical Data Analysis Software and Systems – V*, A.S.P. Conference Series, (eds) Jacoby, G. H. & Barnes, J., **101**, 80.
 Gunn, J. E., Gott, J. R. 1972, *Astrophys. J.*, **176**, 1.
 Haynes, M. P., Giovanelli, R. 1984, *Astron. J.*, **89**, 758.
 Haynes, M. P., Giovanelli, R. 1986, *Astrophys. J.*, **306**, 466.
 Hubble, E., Humason, M. L. 1931, *Astrophys. J.*, **74**, 43.
 Jensen, J. B., Tonry, J. L., Luppino, G. A. 1998, *Astrophys. J.*, **505**, 111.
 Magri, C., Haynes, M. P., Forman, W., Jones, C., Giovanelli, R. 1988, *Astrophys. J.*, **333**, 136.
 Moore, B., Lake, G., Katz, N. 1998, *Astrophys. J.*, **495**, 139.
 Mould J. R., Hughes, S. M. G., Stetson, P. B. *et al.* 2000, *Astrophys. J.*, **528**, 655.
 Mukai, K. 1993, *Legacy 3* (HEASARC Journal), p. 21–31.
 (<http://heasarc.gsfc.nasa.gov/docs/software/tools/pimms.html>)
 Mulchaey, J. S. 2000, *Ann. Rev. Astron. Astrophys.*, **38**, 289.
 Mushotzky, R. F., Smith, B. W. 1980, In *Highlights of astronomy, IAU Proceedings*, **5**, 735.
 Nulsen, P. E. J. 1982, *Mon. Not. R. Astron. Soc.*, **198**, 1007.
 Omar, A. 2004, Ph.D. Thesis, Jawaharlal Nehru University, Delhi.
 Paolillo, M., Andreon, S., Longo, G. *et al.* 2001, *Astron. Astrophys.*, **367**, 59.
 Pellegrini, P. S., da Costa, L. N., Huchra, J. P., Latham, D.W., Willmer, C. N. A. 1990, *Astron. J.*, **99**, 751.
 Poggianti, B. M., Smail, I., Dressler, A. *et al.* 1999, *Astrophys. J.*, **518**, 576.
 Postman, M., Geller, M. J. 1984, *Astrophys. J.*, **281**, 95.
 Rood, H. J., Rothman, V. C. A., Turnrose, B. E. 1970, *Astrophys. J.*, **162**, 411.
 Rupen, M. P. 1999, In *Synthesis imaging in radio astronomy, ASP Conf. Ser.*, **180**, 229.
 Richter, O. G., Sadler, E. M. 1985, *Astron. Astrophys. Suppl.*, **59**, 433.
 Sakai, S., Mould, J. R., Hughes, S. M. G. *et al.* 2000, *Astrophys. J.*, **529**, 698.
 Sandage, A., Tammann, G. A. 1975, *Astrophys. J.*, **196**, 313.
 Spitzer, L. Jr. 1978, In *Physical processes in the Interstellar medium*, (Wiley-interscience publication).
 Swarup, G., Ananthakrishnan, S., Kapahi, V. K. *et al.* 1991, *Curr. Sci.*, **60**, 95.
 Tonry, J. L., Blakeslee, J. P., Ajhar, E. A., Dressler, A. 1997, *Astrophys. J.*, **475**, 399.
 Tonry, J. L. *et al.* 2001, *Astrophys. J.*, **546**, 681.
 Trinchieri, G., Pellegrini, S., Wolter, A., Fabbiano, G., Fiore, F. 2000, *Astron. Astrophys.*, **364**, 53.
 Tully, R. B., Verheijen, M. A. W., Pierce, M. J., Huang, J., Wainscoat, R. J. 1996, *Astron. J.*, **112**, 2471.

- Tully, R. B., Fouque, P. 1985, *Astrophys. J. Suppl.*, **58**, 67.
- Verdes-Montenegro, L., Yun, M. S., Williams, B. A., Huchtmeier, W. K., Del Olmo, A., Perea, J. 2001, *Astron. Astrophys.*, **377**, 812.
- Verheijen, M. A. W., Sancisi, R. 2001, *Astron. Astrophys.*, **370**, 765.
- Warmels, R. H. 1988, *Astron. Astrophys. Suppl.*, **72**, 19.
- Welch, G. A., Chincarini, G., Rood, H. J. 1975, *Astron. J.*, **80**, 77.
- Willmer, C. N. A., Focardi, P., da Costa, L. N., Pellegrini, P. S. 1989, *Astron. J.*, **98**, 1531.

The H_I Content of the Eridanus Group of Galaxies

A. Omar^{*1} & K. S. Dwarakanath²

Raman Research Institute, Sadashivanagar, Bangalore 560 080, India.

¹*e-mail: aomar@upso.ernet.in*

²*e-mail: dwaraka@rri.res.in*

Received 2004 August 23; accepted 2005 March 22

Abstract. The H_I content of galaxies in the Eridanus group is studied using the GMRT observations and the HIPASS data. A significant H_I deficiency up to a factor of 2–3 is observed in galaxies in the high galaxy density regions. The H_I deficiency in galaxies is observed to be directly correlated to the local projected galaxy density, and inversely correlated to the line-of-sight radial velocity. Furthermore, galaxies with larger optical diameters are predominantly in the lower galaxy density regions. It is suggested that the H_I deficiency in Eridanus is due to tidal interactions. In some galaxies, evidences of tidal interactions are seen. An important implication is that significant evolution of galaxies can take place in the group environment. In the hierarchical way of formation of clusters via mergers of groups, a fraction of the observed H_I deficiency in clusters could have originated in groups. The co-existence of S0s and severely H_I deficient galaxies in the Eridanus group suggests that tidal interaction is likely to be an effective mechanism for transforming spirals to S0s.

Key words. Galaxy: evolution—galaxies: groups, clusters—individual: Eridanus—radio lines: H_I 21cm-line.

1. Introduction

Spiral galaxies in the cores of clusters are known to be H_I deficient compared to their field counterparts (Davies & Lewis 1973; Giovanelli & Haynes 1985; Cayatte *et al.* 1990; Bravo-Alfaro *et al.* 2000; Solanes *et al.* 2001). Several gas-removal mechanisms have been proposed to explain the H_I deficiency in cluster galaxies. There are convincing results from both the simulations and the observations that ram-pressure stripping (cf. Gunn & Gott 1972) is effective in galaxies which have crossed the high intra cluster medium (ICM) density region near the core of the cluster (Vollmer *et al.* 2001, van Gorkom 2003). “Galaxy harassment” can also affect outer regions of the disk as a result of repetitive fast encounters of galaxies in clusters (Moore *et al.* 1998). There can be other scenarios where galaxies can become gas deficient, e.g., thermal evaporation and viscous stripping (Cowie & Songaila 1977; Nulsen 1982; Sarazin 1988), and

^{*}Present address: ARIES, Manora peak, Nainital 263 129, Uttarakhand, India.

“galaxy starvation” where hot gas in the halos of galaxies is stripped. It is believed that halos contain a reservoir of hot gas, which sustains star formation in galaxies over their present ages (Larson *et al.* 1980).

Often one or more processes have been shown to be working in individual cases. There are no strong arguments for any of these processes to be globally effective in clusters. Cayatte *et al.* (1990) showed that the H_I deficiency in Virgo galaxies can be understood by a combination of ram-pressure stripping and transport processes. However, there are several inconsistencies. Magri *et al.* (1988) showed that no single gas-removal process can be justified consistently in any cluster. Further uncertainties arise since some of the parameters driving these mechanisms are not known well, e.g., thermal conductivity of the ICM, amount of hot gas in halos, etc. It is also not clear that all severely H_I deficient galaxies have crossed the core as required for ram-pressure to be effective. Contrary to what is expected from ram-pressure stripping, the low mass spirals and dwarfs are indistinguishable from the massive spirals in terms of H_I deficiency (Hoffman *et al.* 1988). Valluri & Jog (1991) observed in Virgo and some other rich clusters that galaxies with medium to large optical sizes tend to be more severely H_I deficient compared to smaller galaxies in terms of both the fractional number and the amount of gas lost. This behaviour is contrary to that expected from ram-pressure stripping or transport processes, however, consistent with that expected if tidal interactions were responsible for the gas deficiency. The exact mechanism responsible for H_I deficiency in clusters is still uncertain. These difficulties have led one to speculate that cluster galaxies were perhaps H_I deficient even before they fell into the cluster. Such a speculation is motivated by the hierarchical theory of structure formation where clusters build up via mergers of small groups. Groups of galaxies therefore provide an opportunity to trace early evolution of galaxies.

Several clusters have been imaged in H_I. However, only limited H_I data exist for large groups. Previous studies on groups were mainly aimed at Hickson Compact Groups (HCGs), which usually have less than 10 galaxies packed in a small volume. For instance, Verdes-Montenegro *et al.* (2001) imaged several HCGs in H_I and found a significant H_I deficiency in the galaxies. To our knowledge, the only large group studied in H_I is Ursa-Major, which is rich in spirals, and has a few S0s and no ellipticals. Verheijen (2001) showed that there is no H_I deficiency in the Ursa-Major galaxies. The environment in the Ursa-Major group is similar to that in the field. The H_I data for groups in which the environment is intermediate between field and cluster is lacking. Here, we present an H_I study of the Eridanus group which appears to be an intermediate system between a loose group like Ursa-major and a cluster like Fornax or Virgo. The properties of the Eridanus group are described in detail in Omar & Dwarakanath (2005a; hereafter paper-I). The Eridanus group has a significant population of S0s (paper-I). The origin of S0s has been the subject of much debate. Usually the population of S0s is enhanced in clusters where galaxy density is high. There are two hypotheses for the formation of S0s, one is “Nature” where it is believed that S0s were formed as such, and the other is “Nurture” (evolution) according to which these galaxies are transformed spirals. The presence of enhanced population of S0s in the Eridanus group indicates that significant evolution of galaxies has perhaps already occurred in the group. In the present study, H_I content of galaxies in the Eridanus group is analysed. The aim is to identify the galaxy evolution processes in the group environment. Both the GMRT data and the HIPASS (HI Parkes All Sky Survey) data are used. The details of the GMRT observations, data reduction and analyses are presented in paper-I and Omar

(2004). Some new results based on follow-up VLA observations are also presented here.

2. The Eridanus group

The Eridanus group was identified as a moderate size cluster in a large scale filamentary structure near $cz \sim 1500 \text{ km s}^{-1}$ in the Southern Sky Redshift Survey (SSRS; da Costa *et al.* 1988). This filamentary structure, which is the most prominent in the southern sky, extends for more than 20 Mpc. The Fornax cluster and the Dorado group of galaxies are also part of this structure. Eridanus has ~ 200 galaxies distributed over ~ 10 Mpc region. The properties of the group are described in detail in paper-I. The distance to the group is estimated as $\sim 23 \pm 2$ Mpc. The group appears to be made of different sub-groups which have different morphological mix. One of the sub-groups, NGC 1407 (cf. Willmer *et al.* 1989), has a population mix of (E+S0s) and (Sp+Irrs) in the ratios of 70% and 30% respectively, which is quite similar to those found in clusters. The overall population mix in the Eridanus group is 30% (E + S0) and 70% (Sp + Irr). These sub-groups often have their brightest member as an elliptical or an S0. The brightest member in the entire group is a luminous ($L_B \sim 4 \times 10^{10} L_\odot$) elliptical galaxy with diffuse X-ray emission ($L_{X; 0.1-2.0 \text{ keV}} \sim 2 \times 10^{41} \text{ erg s}^{-1}$) surrounding it. Diffuse X-ray emission ($L_{X; 0.1-2.0 \text{ keV}} \sim 7 \times 10^{40} \text{ erg s}^{-1}$) is also seen around another elliptical galaxy NGC 1395 which belongs to another sub-group. There is no appreciable difference in the velocities over which the early types and the late types are distributed. This is contrary to that seen in Virgo, Coma and several nearby Abell clusters where spirals have much flatter velocity distribution while E/S0s have nearly a Gaussian distribution in velocity (Binggeli *et al.* 1987; Colless & Dunn 1996; Biviano *et al.* 2002).

3. The H_I content

A total of 57 galaxies in the Eridanus group were observed with the GMRT in the H_I 21 cm-line. The details of the observations and the data analyses are described in paper-I. The H_I detections were made for 31 galaxies. It was noticed that the H_I flux densities of some large (dia. $> 6'$) galaxies were underestimated by the GMRT observations presumably due to inadequate sampling of the short (u, v) spacings. In the present study, the H_I masses for such galaxies were replaced by those obtained from the HIPASS data (Meyer *et al.* 2004). In addition, HIPASS data were used for galaxies in the Eridanus region not observed by the GMRT. The final sample consisted of a total of 63 H_I detected galaxies of different morphological types. The H_I sample is described in Appendix A.

The H_I masses of galaxies in the field environment are observed to be correlated with their Hubble types and optical diameters D_{opt} (e.g., Haynes & Giovanelli 1984, hereafter HG84). The H_I deficiency (cf. HG84) for a galaxy of a given type can be estimated by comparing $\log(M_{\text{H}_I}/D_{\text{opt}}^2)$ of the galaxy with that observed for field galaxies. In the present study, the ratio $\log(M_{\text{H}_I}/D_{\text{opt}}^2)$ for each Eridanus galaxy is compared with the mean value of the ratio $\log(M_{\text{H}_I}/D_{\text{opt}}^2)$ obtained by HG84 for isolated galaxies of similar types. A significant positive difference between the two ratios indicates an H_I deficiency (def. = $\langle \log(M_{\text{H}_I}/D_{\text{opt}}^2) \rangle_{\text{field}} - \log(M_{\text{H}_I}/D_{\text{opt}}^2)$). It should be noted that this deficiency parameter is distance independent. The optical

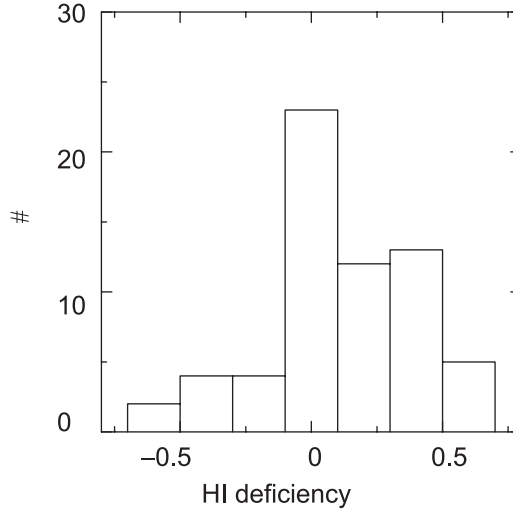


Figure 1. Histogram of the H I deficiency in the Eridanus galaxies.

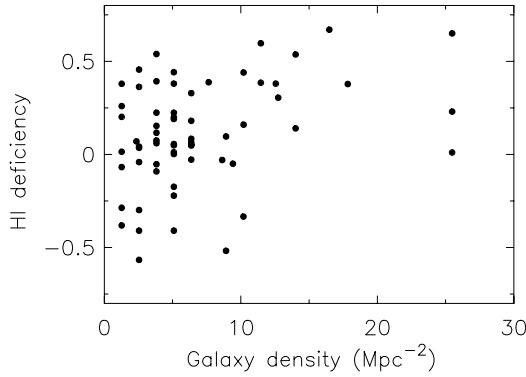


Figure 2. H I deficiency vs. the local projected galaxy density. The Spearman Rank-Order Correlation Coefficient test shows that the correlation is significant at $> 99\%$.

diameters of galaxies used in HG84 were from the Upsala General Catalog (UGC). The optical diameters of galaxies in the Eridanus group are from the Third Reference Catalog of Galaxies (RC3; de Vaucouleurs *et al.* 1991). The optical diameters in RC3 are at 25 mag arc sec⁻² in the B-band. To convert the RC3 diameters or D_{25} to D_{opt} consistent with the UGC diameters, the conversion relation obtained by Paturel *et al.* (1991) was used. This relation predicts that the D_{opt} (UGC) is about 1.09 times the D_{25} .

A histogram of the H I deficiency for the Eridanus galaxies is plotted in Fig. 1. The H I deficiency is independent of the morphological type of the galaxies. It can be seen that although the distribution peaks at zero deficiency, there are more galaxies with positive differences. Some H I rich galaxies (def. < -0.5) are also seen in Fig. 1. These turn out to be interacting pairs, and hence the H I masses are likely to be overestimated.

Figure 2 shows the H I deficiency plotted against the local projected galaxy density. The projected galaxy density is estimated within a circular region of diameter 1.0 Mpc.

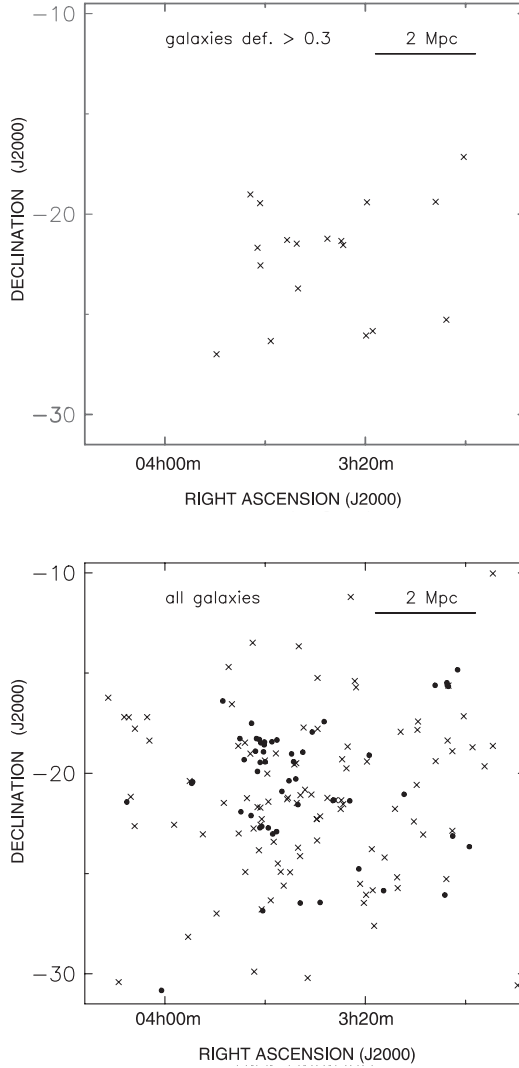


Figure 3. (Lower panel) All galaxies in the Eridanus group. The early type galaxies (E + S0) are marked as filled circles and late type galaxies (Sp + Irr) are marked as crosses. (Upper panel) Galaxies with $H\text{I}$ deficiency greater than 0.3. It can be seen that severely $H\text{I}$ deficient galaxies and early type galaxies are confined to the regions of higher galaxy densities.

It can be seen that in higher galaxy density ($>10 \text{ Mpc}^{-2}$) regions, a majority of galaxies are $H\text{I}$ deficient while in the lower galaxy density regions both normal and deficient galaxies are present. Galaxies are $H\text{I}$ deficient up to a factor of 2–3 ($\sim 0.3\text{--}0.5$ in log units) in these plots. The correlation between the projected galaxy density and $H\text{I}$ deficiency can also be seen in Fig. 3 where the locations of all identified group members (lower panel) and the locations of galaxies deficient by more than a factor of two are plotted (top panel). It is evident that severely $H\text{I}$ deficient galaxies and the early type galaxies are confined to the regions of higher galaxy densities. The $H\text{I}$ deficiency also shows a strong inverse correlation with the line-of-sight radial velocities (w.r.t. the

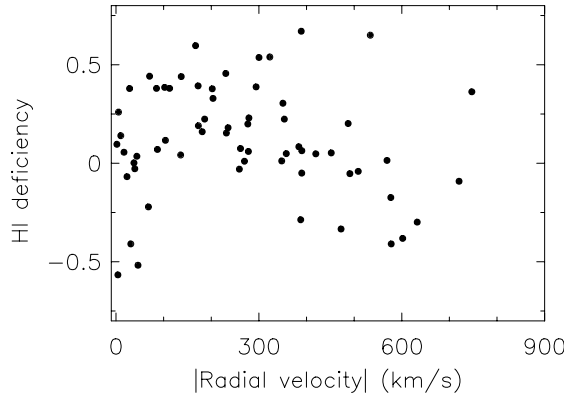


Figure 4. H I deficiency is plotted against the line-of-sight radial velocities (w.r.t. the systemic velocity of the group) of galaxies in the group. The Spearman Rank-Order Correlation Coefficient test shows that the correlation is significant at $> 99.9\%$ for velocities $> 100 \text{ km s}^{-1}$.

systemic velocity of the group) of galaxies in the group (Fig. 4). All of this information is used in the next section to identify the gas-removal mechanism active in the group.

4. The gas-removal processes

Several gas-removal mechanisms have been discussed in the literature to explain the H I deficiency in cluster galaxies. These mechanisms are ram-pressure stripping, transport processes (thermal conduction, and viscous and turbulent stripping), galaxy harassment (tidal interactions), and galaxy strangulation, etc. All these mechanisms are expected to show some correlation of the H I deficiency with the properties of the cluster-environment and of the galaxies. For instance, ram-pressure stripping will be more effective for galaxies with higher radial velocities in the group. The trend observed in Fig. 4 is opposite to that expected if ram-pressure stripping were globally effective in the Eridanus group. It can be shown that for galaxies in the Eridanus group the ram-pressure is one to two orders of magnitude lower than that in the cores of clusters (Omar 2004). This implies that ram-pressure stripping is of little importance in the Eridanus group.

The direct correlation of deficiency with the local projected galaxy density, and the inverse correlation with the line-of-sight radial velocity suggests that the H I deficiency in Eridanus galaxies is due to tidal interactions. Galaxies in higher galaxy density regions will have a higher probability of tidal encounters. Therefore, a direct correlation of deficiency with the local projected galaxy density is expected. Further, the perturbation due to tidal interactions is expected to be maximum for slow encounters. In the Eridanus group where the velocity distribution of galaxies is peaked near the mean velocity of the group (paper-I) and falls off nearly as a Gaussian at higher relative velocities, galaxies having nearly zero radial velocities in the group will have a higher probability of interacting with a companion having a lower velocity difference. Therefore, the inverse correlation of H I deficiency with the line-of-sight radial velocity is qualitatively understood. The presence of a wider distribution of deficiency near the zero radial velocity is likely to be due to projection effects. Galaxies with higher radial velocities but moving nearly perpendicular to the line-of-sight will have almost

zero line-of-sight radial velocities. Therefore, some discordant points are expected near zero velocity in Fig. 4.

Tidal forces will affect both the gas and the stars in galaxies. In contrast, ram-pressure affects only the gas. If the H_I deficiency in the Eridanus group is indeed due to tidal interactions, some observational signatures of the same should be seen in both the stellar and the H_I disks. The tidal interactions often produce gaseous and stellar tidal tails extending to large distances in the IGM. Some of the gas and the stars will be lost from the galaxy to the IGM in this process. It will be difficult to detect this low column density tidal debris in the IGM except for those associated with recent events where the column densities could still be detectable. However, repeated tidal encounters in the high galaxy density regions will shrink the optical sizes of galaxies. Figure 5 indicates that the Eridanus galaxies with larger optical sizes are predominantly in the lower galaxy density regions. This trend is further indicative of the scenario of tidal interactions being effective in the Eridanus group.

It is worthwhile to discuss an important effect while estimating H_I deficiency using the $M_{\text{H I}} / D_{\text{opt}}^2$ parameter. Since both D_{opt} and $M_{\text{H I}}$ are reduced as a result of tidal encounters, the H_I gas loss inferred from this deficiency parameter will be a lower limit. In the absence of detailed simulations of repetitive tidal encounters in a group environment, such effects are hard to quantify.

5. The morphological peculiarities of the Eridanus galaxies

Some galaxies in the Eridanus group show tidal tails and other peculiarities like H_I extending out of the disk, H_I warps, asymmetric H_I disks, shrunken or fragmented H_I disks, kinematical or H_I lopsidedness, etc. Tidal interactions can produce long tails of gas and stars, and deformations in the disks of galaxies. Some representative examples of these peculiarities are shown in Figs. 6 and 7, and are discussed below.

5.1 Shrunken H_I disks

ESO 549-G 002 and NGC 1422 show shrunken H_I disks. Both of these galaxies are H_I deficient, and are in a region with a galaxy density $\sim 20 \text{ Mpc}^{-2}$. The H_I deficiencies

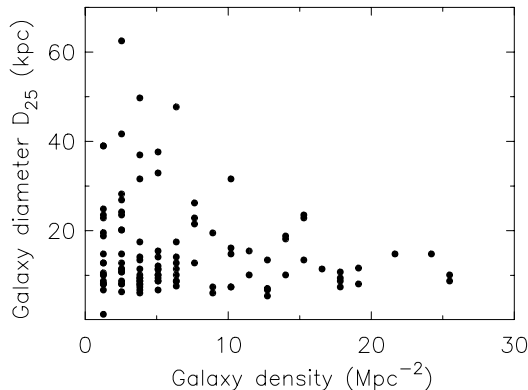


Figure 5. The optical disk diameters plotted against the local projected galaxy density.

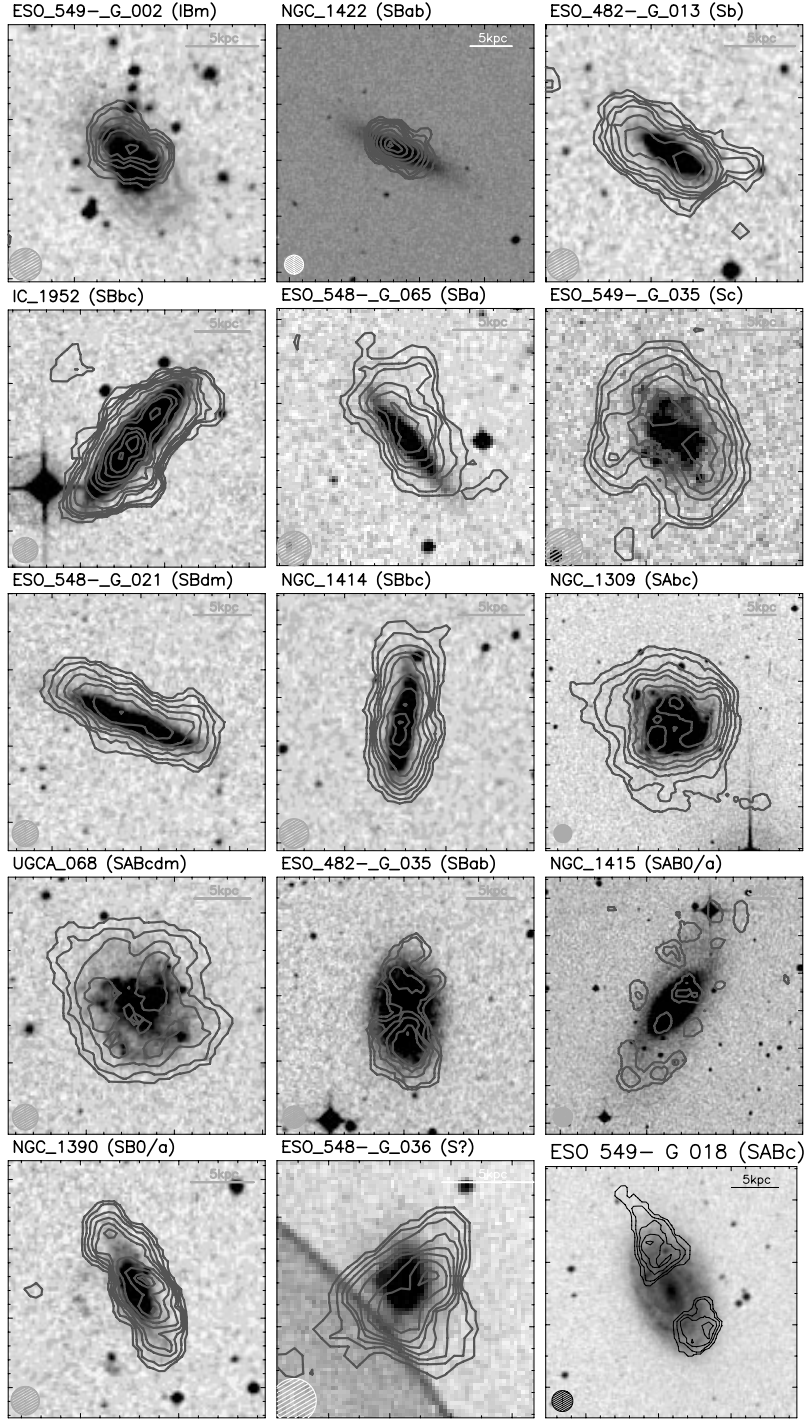


Figure 6. GMRT H I column density contours overlaid upon the optical DSS gray-scale images of Eridanus galaxies which show peculiar H I morphologies (see section 5 for details). The contour levels increase in units of $N_{\text{H I}} = 2 \times 10^{20} \text{ cm}^{-2}$. The first contour is at $N_{\text{H I}} = 10^{20} \text{ cm}^{-2}$.

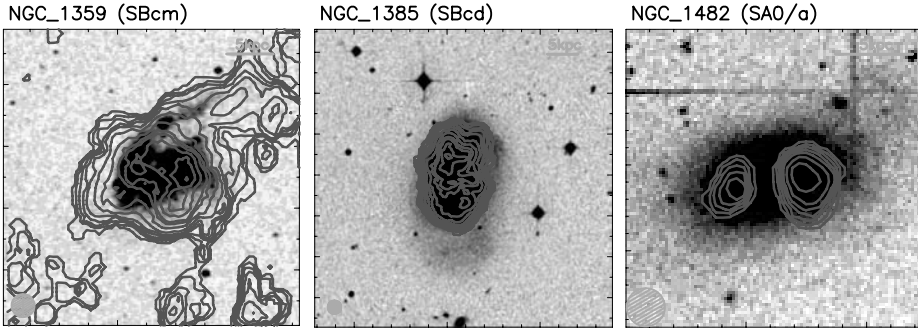


Figure 7. GMRT H_I column density contours overlaid upon the optical DSS gray-scale images of Eridanus galaxies which have clearly visible tidal tails in optical.

for ESO 549-G 002 and NGC 1422 are 0.67 and 0.44 respectively. ESO 549-G 002 has a faint stellar envelop in the outer region and has an irregular optical morphology in the inner region. NGC 1422 is an edge-on galaxy with a prominent dust lane. It has an H_I morphology similar to that seen in ram-pressure stripped galaxies in clusters where gas from the outer regions is preferentially removed. As it has been argued that the ram-pressure is not very effective in the Eridanus group, it indicates that the highly shrunk H_I disk in NGC 1422 is due to some other reasons. Moreover, no tidal features are seen either in optical or in H_I . The origin of the shrunk H_I disk in NGC 1422 therefore remains elusive.

5.2 Extra-planar gas and warps

Some galaxies in the Eridanus group show H_I extending out of the disk (see ESO 482-G 013, IC 1952, ESO 548-G 065, ESO 549-G 035 in Fig. 6). This extra-planar gas is seen in the form of wisps and small plumes of gas. It appears that this phenomena is often seen in the edge-on galaxies since the extra-planar gas will be easier to detect in these systems. It may be possible that other galaxies also have such features. The extra-planar gas might be quite common in galaxies. H_I warps can be noticed in ESO 548-G 021 and NGC 1414.

5.3 Asymmetric H_I disks

NGC 1309 and UGCA 068 have asymmetric H_I disks. These two galaxies have normal H_I content. One side of the galaxy appears more diffuse than the opposite side. These two galaxies also show kinematical asymmetries (paper-I). NGC 1309 has a warp in the outer region. UGCA 068 shows asymmetry in the rotation curve. We speculate that such features could be due to retrograde tidal encounters. The retrograde encounter described by Toomre & Toomre (1972) does not pull out stars (the gas will respond in the same way) but can cause both morphological and kinematical asymmetries.

5.4 Peculiar H_I disks

The H_I disks of ESO 482-G 035 and NGC 1415 are seen to be very peculiar. Both are early type disk galaxies, and it is not common to see fully developed H_I disks in

such galaxies. It is likely that the detected H_I is in a ring rather than being in a fully developed disk. NGC 1415 is an S0/a galaxy with a faint optical ring in the outer regions where most of the H_I is seen in isolated clouds. The inclination of the ring is mis-aligned with the inner disk by more than 20°.

The H_I disk of NGC 1390 is bent in an arc shape. Some diffuse nebulosity is visible towards the west of this galaxy.

5.5 Polar ring galaxy

ESO 548-G 036 (S?) has the position angle of the H_I disk inferred from H_I kinematics almost normal to the position angle of the optical isophotes. The optical body resembles an S0 galaxy with a dust lane normal to the major axis of the main body. It is suggested that this is a polar ring galaxy. Polar rings are believed to be due to recent accretion of gas from tidal encounters. IC 1953, an H_I rich galaxy at a projected separation of ~ 50 kpc, is likely to be the companion.

5.6 Tidal tails

Either gaseous or stellar tidal features are seen in NGC 1359, NGC 1385, and NGC 1482 (Fig. 7). The H_I tidal tail and some isolated H_I features in the vicinity can be seen in NGC 1359. Both NGC 1385 and NGC 1482 are far-infrared luminous galaxies, and undergoing intense star forming activities. NGC 1385 has highly asymmetric diffuse stellar envelop, however no gaseous tidal tail is seen. NGC 1482 (S0/a) appears to have an H_I ring coincident with the dust ring seen in the optical image. The apparent central H_I hole in NGC 1482 is due to H_I absorption against the radio emission. Stellar tidal features can be seen in NGC 1482. Follow-up VLA H_I observations on these galaxies were carried out aimed at detecting any low column density tidal debris. The details are given in the next section.

6. Follow-up VLA observations

Although several peculiarities are seen in both the optical and the GMRT H_I images of Eridanus galaxies, no tidal debris in H_I were detected in the GMRT images. It is expected that tidal debris will have low column density gas which could have been missed in the GMRT images due to its limited sensitivity to the extended emission. Follow-up VLA observations were, therefore, carried out in its D-configuration which is most sensitive to the extended low column density emission. The observations were carried out on galaxies having one or more close neighbours or on galaxies which showed tidal features in optical. The details of the observations are given in Table 1. The observations were carried out in the D-north-C (DnC) hybrid configuration which gives a nearly symmetric synthesised beam for sources in the southern hemisphere. The data were analysed following the standard procedures using AIPS (Astronomical Image Processing System) developed by the National Radio Astronomy Observatory. The flux density scale is based on the standard VLA calibrator 0137+331. Observations were carried out in two polarizations. Fields 2 and 3 (cf. Table 1) were observed in the 4IF correlator mode at two different centre frequencies in the two IFs each of bandwidth 3.125 MHz with sufficient overlap so that the total usable bandwidth after stitching the two IFs was ~ 5 MHz. Other fields were observed in the 2IF correlator

Table 1. VLA D-configuration observations.

#	Obs. date (dd-mm-yy)	Field centre ($\alpha^{h,m,s}$, J2000) ($\delta^{d,m,s}$, J2000)			Galaxies	rms (mJy/bm)	$\theta_a \times \theta_b$, PA ($'' \times ''$, $^\circ$)
1	30-05-04	03	54	49.0	NGC 1482	1.4	74×57 , 15.6
		−20	26	14.0	NGC 1481		
					ESO 549-G 035		
2	05-06-04	03	40	31.5	ESO 548-G 072	1.0	67×58 , 04.5
		−19	25	00.0	ESO 548-G 065		
					ESO 548-G 064		
3	11-06-04	03	41	12.5	NGC 1415	1.0	76×56 , 15.1
		−22	34	30.0	APM 482 + 009-132		
					ESO 482-G 031		
					NGC 1416		
4	12-06-04	03	33	34.8	ESO 548-G 036	1.2	75×56 , 18.6
		−21	31	23.0	IC 1953		
5	13-06-04	03	37	28.3	NGC 1385	1.9	71×57 , −8.6
		−24	30	05.0			
6	13-06-04	03	34	00.0	NGC 1359	1.9	84×55 , 27.0
		−19	28	36.0	ESO 548-G 043		
					ESO 548-G 044		

Notes – (a) The bandwidth of each IF was 3.125 MHz. (b) Fields 1, 4, 5 and 6 were observed in 2 IFs with 128 channels in each IF. Fields 2 and 3 were observed in 4 IFs with 64 channels in each IF. The total integration time was ~ 3.2 h for the fields 1, 2, 3, and 4 and ~ 1.6 h for fields 5 and 6.

mode with 3.125 MHz of bandwidth. The velocity resolution was $\sim 10.4 \text{ km s}^{-1}$ for fields 2 and 3, and $\sim 5.2 \text{ km s}^{-1}$ for the other fields.

The $H\text{I}$ emission was searched by eye in the channel images. The $H\text{I}$ moment maps were constructed using the AIPS task MOMNT. Field 1 (NGC 1482; Fig. 8) and field 6 (NGC 1359; Fig. 9) show extended $H\text{I}$ tails previously undetected in the GMRT images. The $H\text{I}$ images of other galaxies are almost identical to those obtained from the GMRT. Both NGC 1482 and NGC 1359 are in the sub-groups of which they are the brightest members. Both galaxies show tidal features in their optical images. It appears that NGC 1482 has interacted with NGC 1481. NGC 1359 has two nearest neighbours, but $H\text{I}$ streamers do not connect them. It appears that NGC 1359 has undergone multiple interactions as there are two different $H\text{I}$ streamers or tails, one towards the north-east and the other towards the west winding towards the south. An understanding of the $H\text{I}$ morphologies in these galaxies requires detailed N-body simulations. Nevertheless, these $H\text{I}$ detections indicate that tidal interactions are effective in the Eridanus group.

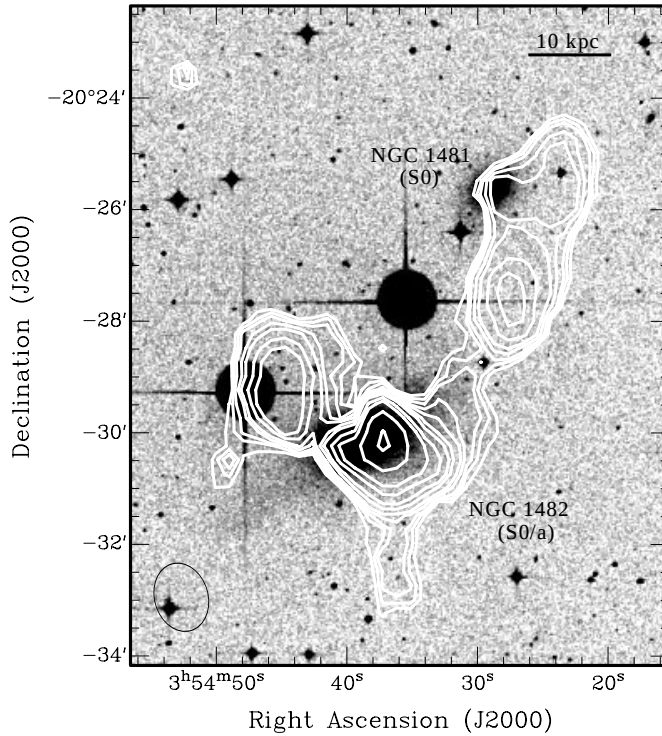


Figure 8. Contours of the VLA H I image of NGC 1482 overlaid upon the optical image from DSS. The contours are at $N_{\text{H I}} = 1, 1.5, 2, 3, 4, 6, 8, 12, 16, 24, 32 \times 10^{19} \text{ cm}^{-2}$. Faint stellar streamers can be seen around NGC 1482.

7. Discussion

Galaxies in the Eridanus group are H I deficient up to a factor of 2–3 compared to their field counterparts. The direct correlation of the H I deficiency with the local galaxy density, and the inverse correlation of the H I deficiency with the line-of-sight radial velocities of galaxies suggest that the deficiency is due to tidal interactions. Since the H I deficiency of the Eridanus galaxies is observed in all types of galaxies, the correlation of the H I deficiency with the local galaxy density cannot be just a manifestation of the density-morphology relation. Although Eridanus appears as a loose group, it has significant sub-grouping. These sub-groups have regions of higher galaxy densities where H I deficient galaxies are seen. A galaxy moving with a radial velocity of $\sim 240 \text{ km s}^{-1}$ (velocity dispersion of the galaxies in the group) can cross a linear distance of $\sim 1 \text{ Mpc}$ (typical extents of the sub-groups) in $\sim 5 \text{ Gyr}$ (typical ages of the galaxies). The higher galaxy density and the relatively short crossing time in the sub-groups enhance the chances of encounters between the galaxies. Therefore, it appears that the sub-grouping is playing an important role in producing the H I deficiency. Further, the inverse correlation of deficiency with the radial velocity indicates that it is necessary for a galaxy to interact with another galaxy with a smaller velocity difference to make the tidal encounters effective.

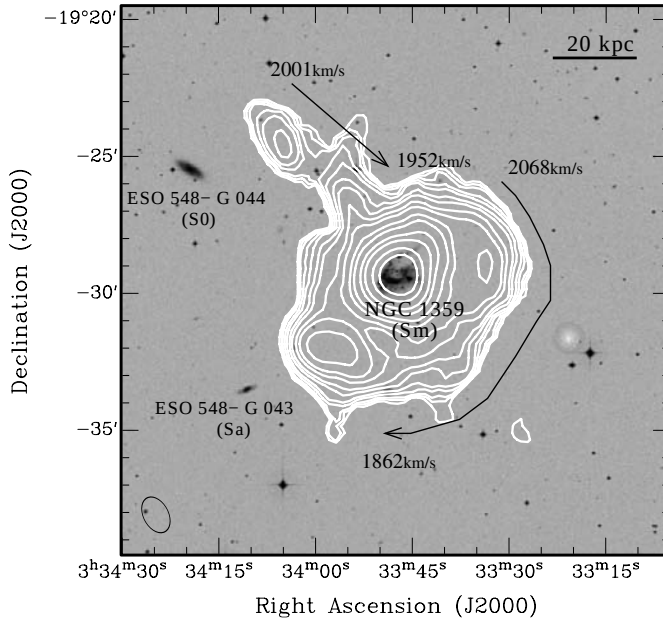


Figure 9. Contours of the VLA $H\text{I}$ image of NGC 1359 overlaid upon the optical image from DSS. The contours are at $N_{H\text{I}} = 1, 1.5, 2, 3, 4, 6, 8, 12, 16, 24, 32, 48, 64, 96, 128 \times 10^{19} \text{cm}^{-2}$. Stellar tidal tail can be seen in NGC 1359. The approximate directions of the velocity gradients in the two tidal features are marked.

The observed $H\text{I}$ deficiency in a loose group like Eridanus has some interesting implications for the evolution of galaxies. Galaxies in clusters are known to be $H\text{I}$ deficient up to a factor of 10. However, the reasons for such large deficiencies are not completely understood. If clusters form as mergers of groups as expected in a hierarchical Universe, a fraction of the $H\text{I}$ deficiency seen in cluster galaxies might have originated in the group environment. In other words, clusters are built with galaxies, some of them are already $H\text{I}$ deficient. This implication is particularly important for simulations which try to match the observed $H\text{I}$ deficiency in cluster galaxies with what can be produced by ram-pressure stripping in the cluster environment (e.g., Vollmer *et al.* 2001). Tidal interactions will become less effective in removing matter from galaxies in a cluster environment as encounters will be relatively fast. However, increased frequency of encounters in clusters due to higher galaxy density may actually affect the outer regions of galaxies as envisaged by the process of “galaxy harassment”. Therefore, it is possible that tidal interactions play an important role in the evolution of galaxies in both the group and the cluster environments.

The origin of S0s in high galaxy density regions is not understood. Whether S0s are due to some evolutionary processes driven by the environment (“Nurture”) or due to the result of natural processes of galaxy formation (“Nature”) is being debated. Several co-workers (Dressler *et al.* 1997; Poggianti *et al.* 1999; Fasano *et al.* 2000) found that in observations of clusters at intermediate redshifts ($z \sim 0.1 - 0.3$), the fraction of S0s tends to grow at the expense of the spiral population as the redshift decreases. These observations support the “Nurture” scenario. Several mechanisms have been proposed for the transformation of spirals into S0s, e.g., ram-pressure stripping, galaxy

harassment, strangulation. The present observations in the Eridanus group indicate that both the severely H_I deficient galaxies and the S0s are found in the higher galaxy density regions in the Eridanus group. It may be an indication that both the H_I deficient and the S0 galaxies originated through similar processes. Since the ram-pressure is not playing any major role in the Eridanus group, the S0s in Eridanus are produced by some other process. Tidal interactions appear to be a favourable mechanism for transformation of spirals into S0s in the wake of the current understanding of the Eridanus group from this study.

The H_I deficiency observed in the Eridanus galaxies is consistent with the H_I deficiency seen in other low velocity dispersion groups and clusters, viz., the Hickson Compact groups (Verdes-Montenegro *et al.* 2001) and the Fornax cluster (Schroder *et al.* 2001). Mulchaey (2000) observed that groups of galaxies are filled with the metal-enriched intra-group medium with average metallicity ~ 0.3 . While other mechanisms can also enrich the intra-group medium with metals, gas lost from the galaxies via tidal interactions will have some contribution to the metal content of the intra-group medium.

The environment in the Eridanus group is intermediate between that in a loose group like the Ursa-Major and a cluster like the Fornax or the Virgo (paper-I). The Ursa-Major has a few S0s and the H_I contents of spirals are similar to the field spirals. However, the Eridanus group which has a relatively large velocity dispersion, and a relatively large fraction of S0s, shows significant H_I deficiency. Willmer *et al.* (1989) indicated that the Eridanus group is in its early phase of cluster formation. If all these results are combined, it appears that the Eridanus group is indeed in an initial phase of cluster formation where physical conditions in the group are favourable for driving the galaxy evolution.

8. Conclusions

- The galaxies in the Eridanus group are H_I deficient up to a factor of 2–3. The H_I deficiency is inversely correlated to the line-of-sight radial velocity, and directly correlated to the local projected galaxy density.
- The H_I deficiency in the Eridanus group is likely to be due to tidal interactions.
- The optical diameters of galaxies are observed to be reduced in the high galaxy density regions indicating the effectiveness of tidal interactions in the Eridanus group.
- If clusters are built via mergers of groups, a fraction of the H_I deficiency in cluster galaxies might have been produced in the group environment.
- The co-existence of S0s and H_I deficient galaxies in the Eridanus group suggests that tidal interactions may be an effective mechanism for transforming spirals to S0s.
- The environment in the Eridanus group is intermediate between the field and a cluster. Nevertheless, conditions are favourable for driving galaxy evolution in the Eridanus group.

Acknowledgements

We thank Marc Verheijen, Jacqueline van Gorkom, and the referee for useful comments. This research has made use of the HI Parkes All Sky Survey (HIPASS) data and has been benefited by NASA's Astrophysics Data System (ADS) and Extra-galactic Database (NED) services.

Appendix A: The H I detected galaxies**Table 2.**

Galaxy	H.T.	D_{UGC} (kpc)	Velocity (km s ⁻¹)	Proj. density (Mpc ⁻²)	$\log(M_{H I} / M_{\odot})$	Deficiency
NGC 1076	2	13.9	2102	1.3	9.27	-0.38
NGC 1069	3	25.5	1456	2.5	9.47	0.04
NGC 1140	9	12.4	1496	2.5	9.62	-0.57
ESO 546-G 034	9	9.5	1568	5.1	9.05	-0.22
NGC 1163	6	21.2	2247	2.5	9.14	0.36
NGC 1187	7	40.1	1396	3.8	9.88	0.12
NGC 1179	8	35.7	1777	5.1	9.78	0.20
UGCA 050	9	13.9	1731	2.5	8.70	0.46
UGCA 051	9	12.4	1672	3.8	8.66	0.39
ESO 547-G 011	9	9.5	2220	3.8	8.92	-0.09
IC 1898	7	26.3	1327	5.1	9.44	0.19
ESO 547-G 020	9	10.2	1991	3.8	8.94	-0.05
NGC 1255	6	30.6	1686	3.8	9.60	0.22
UGCA 061	9	29.2	1735	6.4	9.62	0.18
NGC 1258	8	9.5	1483	5.1	8.77	0.06
UGCA 063	9	8.8	2077	5.1	8.93	-0.17
NGC 1292	7	21.9	1364	2.5	9.43	0.04
UGCA 064	8	21.1	1794	7.6	9.13	0.39
UGCA 065	9	14.6	1537	5.1	9.20	0.00
NGC 1300	6	45.2	1571	5.1	9.72	0.44
NGC 1302	3	28.4	1704	6.4	9.27	0.33
NGC 1306	5	8.0	1454	8.9	9.16	-0.52
MCG-03-09-027	9	11.7	2009	2.5	9.05	-0.04
NGC 1309	6	16.0	2132	2.5	9.56	-0.30
UGCA 068	9	12.4	1848	5.1	9.05	0.01
NGC 1325	6	34.3	1602	11.5	9.53	0.39
NGC 1325A	6	16.0	1333	11.5	8.66	0.60
ESO 548-G 021	8	14.6	1702	17.8	8.82	0.38
NGC 1345	7	10.9	1531	5.1	9.28	-0.41
UGCA 075	9	13.9	1888	1.3	9.44	-0.29
UGCA 077	9	12.4	1952	6.4	9.00	0.05
ESO 482-G 005	8	12.3	1920	6.4	9.00	0.05
NGC 1357	4	20.4	1987	1.3	9.10	0.20

Table 2. (Continued)

Galaxy	H.T.	D_{UGC} (kpc)	Velocity (km s ⁻¹)	Proj. density (Mpc ⁻²)	$\log(M_{H\text{I}}/M_{\odot})$	Deficiency
IC 1952	6	19.0	1823	3.8	8.87	0.54
IC 1953	8	20.4	1851	12.7	9.19	0.30
NGC 1359	9	17.5	1973	10.2	9.69	-0.33
NGC 1371	3	40.8	1461	6.4	9.94	-0.03
IC 1962	8	19.7	1800	14.0	8.92	0.54
ESO 482-G 013	5	8.2	1854	5.1	8.44	0.22
NGC 1385	8	24.8	1498	8.9	9.56	0.10
NGC 1390	3	10.2	1230	25.5	8.70	0.01
NGC 1398	4	51.8	1388	5.1	9.74	0.38
ESO 482-G 032	9	10.9	1732	3.8	8.79	0.15
NGC 1425	5	42.3	1505	1.3	9.82	0.26
NGC 1421	6	25.5	2069	1.3	9.65	0.01
MCG-03-10-045	9	9.5	1239	3.8	8.75	0.07
UGCA 085	7	24.8	1529	1.3	9.20	0.38
ESO 549-G 035	7	10.2	1794	3.8	9.05	-0.24
IC 2007	7	9.5	1523	1.3	8.81	-0.07
SGC 0401.3-1720	9	12.4	1890	6.4	8.99	0.06
UGCA 087	9	16.8	1884	6.4	9.24	0.08
NGC 1518	8	21.9	922	2.5	9.96	-0.41
UGCA 088	8	13.9	1858	5.1	9.10	0.05
ESO 548-G 049	?	6.7	1510	14.0	8.47	0.14
ESO 549-G 0.35	5	9.4	1778	3.8	8.72	0.06
ESO 548-G 065	1	10.1	1221	25.5	8.46	0.23
NGC 1347	5	7.1	1759	8.6	8.64	-0.03
NGC 1415	1	25.8	1585	12.6	9.05	0.38
NGC 1414	4	11.6	1681	10.2	8.76	0.16
ESO 548-G 072	5	8.7	2034	25.5	8.30	0.65
ESO 482-G 035	2	12.8	1890	9.4	8.64	-0.05
NGC 1422	3	14.8	1637	10.2	8.49	0.44
ESO 549-G 002	9	8.7	1111	16.5	8.27	0.67
ESO 549-G 018	5	17.5	1587	2.4	8.63	0.07

Column 1 – Galaxy name; **Column 2** – Hubble type; **Column 3** – UGC optical diameter; **Column 4** – Systemic velocity; **Column 5** – Projected galaxy density; **Column 6** – H_I mass; **Column 7** – H_I deficiency.

References

- Binggeli, B., Tammann, G. A., Sandage, A. 1987, *Astron. J.*, **94**, 251.
- Biviano, A., Katgert, P., Thomas, T., Adami, C. 2002, *Astron. Astrophys.*, **387**, 8.
- Bravo-Alfaro, H., Cayatte, V., van Gorkom, J. H., Balkowski, C. 2000, *Astron. J.*, **119**, 580.
- Cayatte, V., van Gorkom, J. H., Balkowski, C., Kotanyi, C. 1990, *Astron. J.*, **100**, 604.
- Colless, M., Dunn, A. M. 1996, *Astrophys. J.*, **458**, 435.
- Cowie, L. L., Songaila, A. 1977, *Nature*, **266**, 501.
- da Costa, L. N., Pellegrini, P. S., Sargent, W. L. *et al.* 1988, *Astrophys. J.*, **327**, 544.
- Davies, R. D., Lewis, B. M. 1973, *Mon. Not. R. Astron. Soc.*, **165**, 231.
- de Vaucouleurs *et al.* 1991, Third Reference Catalogue of Bright Galaxies, Springer Verlag, vol. **1–3**.
- Dressler, A., Oemler, A. Jr., Couch, W. J. *et al.* 1997, *Astrophys. J.*, **490**, 577.
- Fasano, G., Poggianti, B. M., Couch, W. J. *et al.* 2000, *Astrophys. J.*, **542**, 673.
- Giovanelli, R., Haynes, M. P. 1985, *Astrophys. J.*, **292**, 404.
- Gunn, J. E., Gott, J. R. 1972, *Astrophys. J.*, **176**, 1.
- Haynes, M. P., Giovanelli, R. 1984, *Astron. J.*, **89**, 758.
- Hoffman, G. L., Helou, G., Salpeter, E. E. 1988, *Astrophys. J.*, **324**, 75.
- Larson, R. B., Tinsley, B. M., Calswell, C. N. 1980, *Astrophys. J.*, **237**, 692.
- Magri, C., Haynes, M. P., Forman, W., Jones, C., Giovanelli, R. 1988, *Astrophys. J.*, **333**, 136.
- Meyer, M. J., Zwaan, M. A., Webster, R. L. *et al.* 2004, *Mon. Not. R. Astron. Soc.*, **350**, 1195.
- Moore, B., Lake, G., Katz, N. 1998, *Astrophys. J.*, **495**, 139.
- Mulchaey, J. S. 2000, *Ann. Rev. Astron. Astrophys.*, **38**, 289.
- Nulsen, P. E. J. 1982, *Mon. Not. R. Astron. Soc.*, **198**, 1007.
- Omar, A. 2004, Ph.D. thesis, Jawaharlal Nehru University, Delhi.
- Omar, A., Dwarakanath, K. S. 2005a (paper-I), *J. Astrophys. Astr.*, this volume, 1.
- Paturel, G., Garcia, A. M., Fouque, P., Buta, R. 1991, *Astron. Astrophys.*, **243**, 319.
- Poggianti, B. M., Smail, I., Dressler, A. *et al.* 1999, *Astrophys. J.*, **518**, 576.
- Sarazin, C. L. 1988, X-ray emission from clusters of galaxies, Cambridge Astrophys. Series, Cambridge University Press.
- Schroder, A., Drinkwater, M. J., Richter, O.-G. 2001, *Astron. Astrophys.*, **376**, 98.
- Solanes, J. M., Manrique, A., Garca-Gmez, C. *et al.* 2001, *Astrophys. J.*, **548**, 97.
- Toomre, A., Toomre, J. 1972, *Astrophys. J.*, **178**, 623.
- Valluri, M., Jog, C. J. 1991, *Astrophys. J.*, **374**, 103.
- van Gorkom, J. H. 2003, In *Clusters of galaxies: Probes of cosmological structure and galaxy formation*, (eds) Mulchaey, J. S., Dressler, A., Oemler, A. (Carnegie Obs. Astrophys. Ser.), **3**.
- Verdes-Montenegro, L., Yun, M. S., Williams, B. A., Huchtmeier, W. K., Del Olmo, A., Perea, J. 2001, *Astron. Astrophys.*, **377**, 812.
- Verheijen, M. A. W. 2001, In *Gas and galaxy evolution*, (ed.) Hibbard, J. E. *et al.*, ASP Conf. Series, **240**, 573.
- Vollmer, B., Cayatte, V., Balkowski, C., Duschl, W. J. 2001, *Astrophys. J.*, **561**, 708.
- Willmer, C. N. A., Focardi, P., da Costa, L. N., Pellegrini, P. S. 1989, *Astron. J.*, **98**, 1531.

Radio Continuum and Far-infrared Emission from the Galaxies in the Eridanus Group

A. Omar^{*1} & K. S. Dwarakanath²

Raman Research Institute, Sadashivanagar, Bangalore 560 080, India.

¹e-mail: aomar@upso.ernet.in

²e-mail: dwaraka@rri.res.in

Received 2004 August 23; accepted 2005 March 22

Abstract. The Eridanus galaxies follow the well-known radio–FIR correlation. The majority (70%) of these galaxies have their star formation rates below that of the Milky Way. The galaxies that have a significant excess of radio emission are identified as low luminosity AGNs based on their radio morphologies obtained from the GMRT observations. There are no powerful AGNs ($L_{20\text{ cm}} > 10^{23} \text{ W Hz}^{-1}$) in the group. The two most far-infrared and radio luminous galaxies in the group have optical and H α morphologies suggestive of recent tidal interactions. The Eridanus group also has two far-infrared luminous but radio-deficient galaxies. It is believed that these galaxies are observed within a few Myr of the onset of an intense star formation episode after being quiescent for at least a 100 Myr. The upper end of the radio luminosity distribution of the Eridanus galaxies ($L_{20\text{ cm}} \sim 10^{22} \text{ W Hz}^{-1}$) is consistent with that of the field galaxies, other groups, and late-type galaxies in nearby clusters.

Key words. Galaxy: radio continuum, radio–FIR correlation—galaxies: groups—individual: Eridanus.

1. Introduction

There is a well-known correlation between the non-thermal radio emission and the thermal far-infrared (FIR) emission from star forming galaxies (van der Kruit 1973; Condon 1992; Yun *et al.* 2001). The radio–FIR correlation has been observed to be extremely tight ($\sigma \sim 0.26$ dex) over a wide range (10^9 – $10^{12} L_\odot$) of FIR luminosities (Yun *et al.* 2001). The origin and the tightness of the radio–FIR correlation is not completely understood. The radio–FIR correlation is linked to the star formation activities in galaxies, and is not displayed by Active Galactic Nuclei (AGNs). According to one scenario, ultra-violet photons from massive stars ($M > 8M_\odot$) heat the surrounding dust which emits in the far-infrared, and the same population of massive stars undergoes supernova explosions thereby accelerating cosmic electrons responsible for much

^{*}Present address: ARIES, Manora peak, Nainital 263 129, Uttaranchal, India.

of the radio emission from galaxies (Harwit & Pacini 1975). The presence or absence of the correlation therefore can be used to make a distinction between the star formation and the AGN related radio continuum emission in galaxies.

The radio–FIR correlation has also been used to study the effects of environment on the evolution of radio sources in galaxy clusters and groups (e.g., Menon 1995; Scodreggio & Gavazzi 1993; Gavazzi & Boselli 1999; Miller & Owen 2001; Reddy & Yun 2004). Studies indicate that normal star forming galaxies in the local Universe seldom have their 20 cm spectral luminosities exceeding $10^{23} \text{ W Hz}^{-1}$, but AGNs are observed with higher radio luminosities (Condon 1989; Yun *et al.* 2001). Studies in clusters of galaxies indicate that the luminous radio AGNs ($L_{20 \text{ cm}} > 10^{23} \text{ W Hz}^{-1}$) are hosted mostly by early type galaxies in the cores of clusters (e.g., Gavazzi & Boselli 1999; Reddy & Yun 2004). These radio-luminous AGNs are more commonly found in the cluster environment than in the field. These studies indicate that the upper end of the radio luminosity distribution of late type galaxies does not differ much in different environments.

The Eridanus group was recently studied in detail (Omar 2004; Omar & Dwarakanath 2005a & b). This group has ~ 200 identified members. This group was predicted to be in an early phase of the cluster formation where different sub-groups were merging together (Willmer *et al.* 1989). Although the entire group is not dynamically relaxed, the galaxy evolution process seems to be effective in the group. A significant H I deficiency up to a factor of 2–3 was observed in the Eridanus galaxies in the high galaxy density regions (Omar & Dwarakanath 2005b). This group also has a significantly large (30%) population of early types (E + S0s). In this paper, the radio–FIR correlation for the Eridanus galaxies is constructed using the FIR data from the *Infrared Astronomical Satellite* (IRAS; Neugebauer *et al.* 1984) survey and the 20 cm radio data from the *Northern VLA Sky Survey* (NVSS; Condon *et al.* 1998). The higher resolution ($5''$ – $10''$) radio continuum images at 20 cm from the GMRT observations are used to study the nature of the galaxies departing from the radio–FIR correlation. The GMRT radio continuum images of some of the normal star forming galaxies are also presented.

2. The FIR and the radio data

2.1 The IRAS data

The IRAS faint source catalog has a limiting sensitivity of $\sim 0.3 \text{ Jy}$ at $\lambda 60\mu$. The 60μ flux density, $S_{60\mu}$ is converted to the luminosity, $L_{60\mu}$ using the relation:

$$\log \frac{L_{60\mu}}{L_{\odot}} = 6.014 + 2 \log \frac{D}{\text{Mpc}} + \log \frac{S_{60\mu}}{\text{Jy}}. \quad (1)$$

At the distance to the Eridanus group ($D \sim 23 \text{ Mpc}$), the limiting sensitivity of the IRAS observations is $L_{60\mu} \sim 10^8 L_{\odot}$. The total FIR luminosity, L_{FIR} is estimated using the relation (Helou *et al.* 1998):

$$L_{\text{FIR}}(L_{\odot}) = \left(1 + \frac{S_{100\mu}}{2.58 S_{60\mu}} \right) L_{60\mu}. \quad (2)$$

2.2 The NVSS data

The 20 cm flux densities are from the NVSS made at a resolution of $45''$ (~ 5 kpc) with an rms of $0.5 \text{ mJy beam}^{-1}$. The 20 cm spectral luminosity, $L_{20 \text{ cm}}$ is estimated using the relation:

$$\log \frac{L_{20 \text{ cm}}}{\text{W Hz}^{-1}} = 20.08 + 2 \log \frac{D}{\text{Mpc}} + \log \frac{S_{20 \text{ cm}}}{\text{Jy}}. \quad (3)$$

The 5σ sensitivity to the 1.4 GHz spectral luminosity is $\sim 1.6 \times 10^{20} \text{ W Hz}^{-1}$.

2.3 The GMRT observations

A total of 57 galaxies in the Eridanus group were observed by Omar & Dwarakanath (2005a) using the GMRT. The observing strategy was optimized to get uniform distribution of visibilities. Two galaxies were observed alternately for 15–20 min each followed by 5–7 min of observations of the secondary calibrators. This cycle was repeated and a total of 3–4 h of observing time was accumulated on each galaxy. A majority of these observations were carried out over a bandwidth of 8 MHz divided into 128 spectral channels. These observations were primarily aimed at detecting H_I emission from the Eridanus galaxies. No pre-selection of galaxies was made based on their radio or optical properties. Both early type and late type galaxies were observed. The data were corrected for the spectral response of the IF filters. The channels devoid of H_I line signals were averaged to obtain the radio continuum data. The GMRT has a mix of both short and long baselines which make the data sensitive to extended emission and at the same time also make it possible to obtain a high resolution image. The data were analysed following the standard procedures using the *Astronomical Image Processing System* (AIPS) developed by the National Radio Astronomy Observatory (NRAO). The flux density scale of these observations is based on the standard VLA calibrators 3C 48 and 3C 147. Images were CLEANed and self-calibrated to improve the dynamic range. The images presented here are at angular resolutions which emphasize the extended radio continuum emission from the galaxies. The details of the image properties are given in Table 1. The total flux densities are estimated from images at a resolution of $\sim 1'$. The GMRT flux densities are consistent with the NVSS flux densities within 20% for most of the galaxies except for the two weaker ones, viz., IC 1953 and ESO 548-G 036 where the uncertainties are 25% and 30% respectively.

3. The radio and FIR emission from the Eridanus galaxies

A total of 72 Eridanus galaxies are detected in the IRAS survey, out of which 38 galaxies were detected in the NVSS. A total of 7 early type (E + S0s) galaxies are detected in radio, most of them being radio weak ($L_{20 \text{ cm}} \sim 10^{20} \text{ W Hz}^{-1}$). The list of Eridanus galaxies detected in the FIR and in the radio is given in Table 2. The histograms of the radio and the FIR luminosities of the galaxies in the Eridanus group are plotted in Fig. 1. The Eridanus galaxies have their radio spectral luminosities below $\sim 10^{22} \text{ W Hz}^{-1}$. The high end of the radio luminosity of the galaxies in the Hickson compact groups (Menon 1995) and the Ursa-Major group (Verheijen & Sancisi 2001) is consistent with that in the Eridanus galaxies. A majority of the late-type galaxies in nearby clusters

Table 1. GMRT image parameters.

Galaxy	rms (mJy/bm)	Resolution (arc sec)	S _{total} (mJy)	Morph.
NGC 1407	0.50	6 × 6	99 ± 10	Diffuse
NGC 1371	0.22	15 × 15	19.7 ± 2	Linear
NGC 1415	0.12	4 × 4	27 ± 3	Linear
NGC 1482	0.81	8 × 8	280 ± 30	Diffuse
NGC 1385	0.41	15 × 15	180 ± 20	Diffuse
NGC 1377	0.20	15 × 15	< 1 (5σ)	–
IC 1953	0.30	50 × 45	9 ± 2	Diffuse
NGC 1309	0.35	15 × 15	68 ± 7	Diffuse
NGC 1345	0.23	32 × 26	4.5 ± 1	Diffuse
ESO 548-G 036	0.20	16 × 11	9 ± 1	Diffuse

Notes: Images are natural-weighted. The resolution of each image is such that the diffuse emission is emphasized without losing much of the resolution.

also have their radio luminosities below $\sim 10^{22} \text{ W Hz}^{-1}$ (Reddy & Yun 2004). It is interesting to observe that the groups (Ursa-Major, HCG's, and Eridanus) lack in powerful AGNs ($L_{20 \text{ cm}} > 10^{23} \text{ W Hz}^{-1}$), which are more commonly associated with the early-type galaxies in clusters (Reddy & Yun 2004). It can also be noticed from Fig. 1 that the majority ($\sim 70\%$) of the Eridanus galaxies have their star formation rates below that of the Milky Way ($\sim 1 \text{ M}_{\odot} \text{ yr}^{-1}$).

4. The radio–FIR correlation

The radio–FIR correlation for the Eridanus galaxies is shown in Fig. 2. The 5σ upper limits ($\log (L_{20 \text{ cm}}/\text{W Hz}^{-1}) \sim 20.2$) for the non-detections in radio are marked by arrows. The sample size for the Eridanus galaxies is relatively small, and does not span a large range of infrared luminosity. Therefore, any fit is of little statistical importance. The best fit straight line for the radio–FIR correlation obtained by Yun *et al.* (2001) on a large sample of IRAS galaxies is shown in Fig. 2. This straight line corresponds to a slope of 1 and an intercept of 12.07. It can be seen that a majority of the Eridanus galaxies follow the radio–FIR correlation.

Following Condon *et al.* (1991), the q parameter is plotted for the Eridanus galaxies in Fig 3. The q parameter is estimated using the following relation:

$$q = \log \left(\frac{2.58 S_{60\mu} + S_{100\mu}}{2.98 \text{ Jy}} \right) - \log \left(\frac{S_{20 \text{ cm}}}{\text{Jy}} \right). \quad (4)$$

In Fig. 3, the two dotted lines mark three times radio excess (below the mean) and three time FIR excess (above the mean) respectively. The mean value shown by the solid line is at $q = 2.34$ as obtained by Yun *et al.* (2001).

It can be noticed from Figs. 2 and 3 that there are some galaxies which deviate significantly (more than 3 times) from the mean radio–FIR correlation. The rest of the galaxies follow the radio–FIR correlation within an rms of 0.47 dex (Fig. 2). A significant

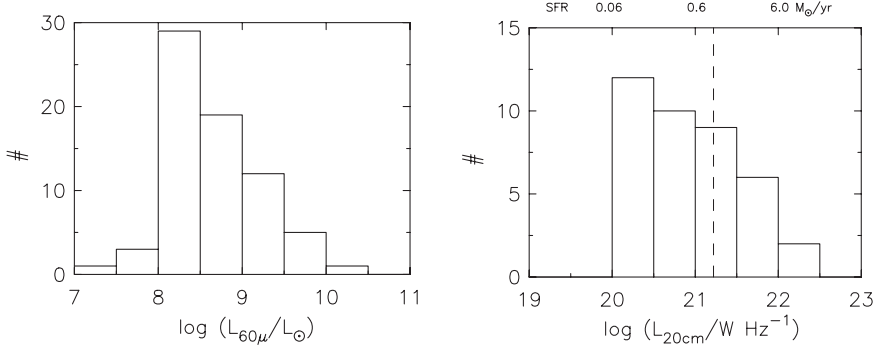


Figure 1. Histograms of radio and FIR luminosities of the Eridanus galaxies. The dashed vertical line in the right hand side panel corresponds to the star formation rate of the Milky Way.

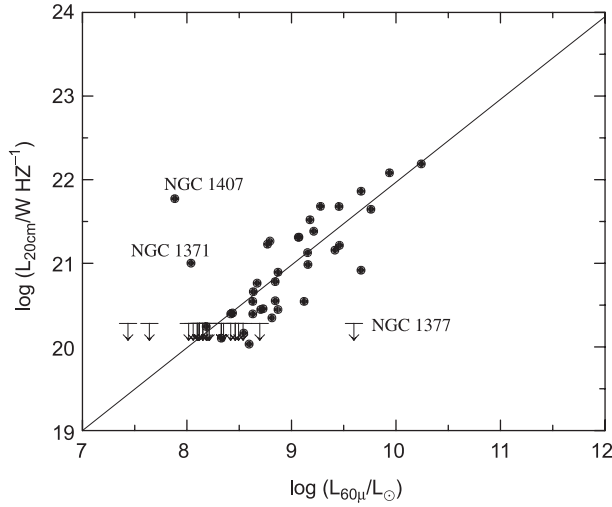


Figure 2. The radio–FIR correlation for the galaxies in the Eridanus group using the data from IRAS and NVSS. The straight line is not a fit to the data. It is the best fit obtained by Yun *et al.* (2001) for an all-sky sample of infrared detected galaxies from IRAS.

deviation in the q value can be noticed at FIR luminosities below $10^9 L_{\odot}$. Such trends are known from earlier studies (e.g., Condon *et al.* 1991; Yun *et al.* 2001). It is believed that this non-linearity is due to the heating of dust from the old stellar population in galaxies (Condon *et al.* 1991). The FIR emission due to this heating is insignificant compared to that due to heating from massive stars in normal star forming galaxies.

5. The radio-excess galaxies

The two Eridanus galaxies, viz., NGC 1407 and NGC 1371 have significant radio excesses. These two galaxies are marked in Fig. 3 (also, see Table 3). The radio

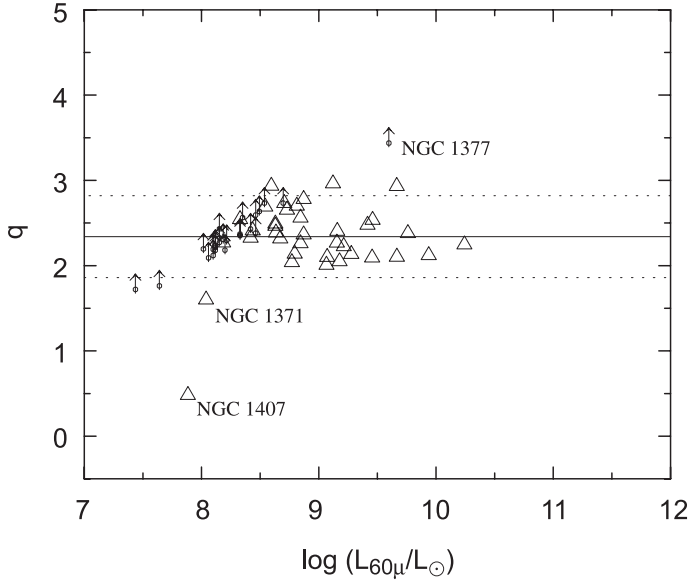


Figure 3. The q parameter for the Eridanus galaxies. The solid line is at $q = 2.34$ which is the mean value obtained by Yun *et al.* (2001), and is not a fit for the present data. The top and the bottom dotted lines are limits for three times FIR excess and three times radio excess from the mean respectively.

emission in excess of that expected from the mean radio–FIR correlation may be due to an AGN in the galaxy. The radio continuum emission from AGNs is often from a nuclear point source or a pair of jets or from both. The GMRT radio continuum contour images of these two galaxies are shown in Fig. 4 overlaid upon their respective optical images from the DSS. It can be noticed that both the galaxies have kpc-scale linear radio structures in their respective centres. Such a feature is indicative of an AGN. A description of these two galaxies is given below.

5.1 NGC 1407 (E0)

NGC 1407 has more than 70 times excess radio emission compared to that expected from the radio–FIR correlation. It is the most optically luminous ($L_B = 4 \times 10^{10} L_\odot$) galaxy in the Eridanus group. It resides in a sub-group having a morphological mix of 70% (E + S0s) and 30% (Sp + Irrs) which is similar to those in galaxy clusters. NGC 1407 has a nuclear X-ray source. Diffuse X-ray emission is also seen in the intra-group medium surrounding the galaxy (Omar & Dwarakanath 2005a). This galaxy was not detected in H I ($M_{\text{H I}}(5\sigma) < 1.2 \times 10^7 M_\odot$). No H I absorption was detected towards the radio source down to a 5σ optical depth limit of 0.05.

The GMRT radio continuum morphology (Fig. 4) shows diffuse emission along a linear structure and a nuclear continuum source coincident with the X-ray nuclear source. The X-ray and radio nucleus, a linear radio structure, and the large radio-excess suggest the presence of an AGN in this galaxy.

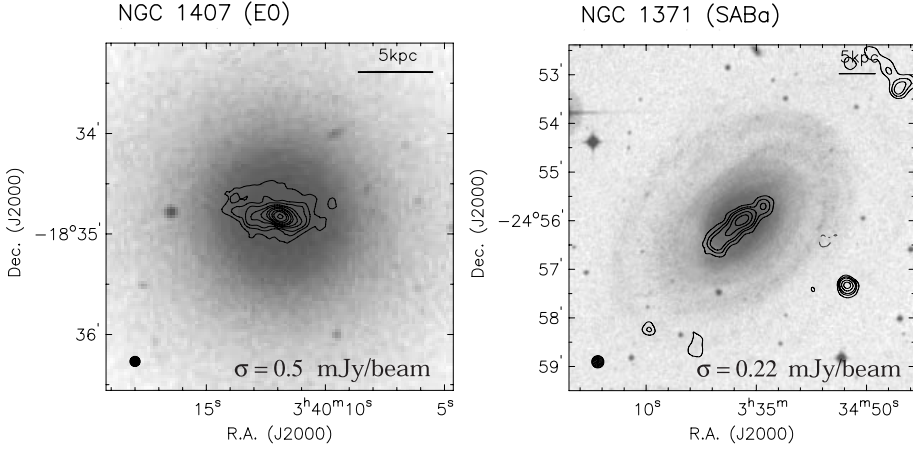


Figure 4. The contours of the radio continuum emission from the radio-excess galaxies in the Eridanus group observed with the GMRT are overlaid upon their respective optical images from the DSS. The contours start at 3 times the rms value (σ , indicated in the map) and increases in steps of 4.5, 6, 9, 12, 18, 24, 36, 48, $72 \times \sigma$.

5.2 NGC 1371 (SABa)

NGC 1371 has more than 5 times excess radio emission compared to that expected from the radio–FIR correlation. It is among one of the largest ($D_{B25} \sim 38$ kpc) galaxies in the Eridanus group. The optical luminosity ($L_B = 2 \times 10^{10} L_\odot$) is dominated by the bulge. It has faint spiral arms in the outer regions. The galaxy was observed in $H\alpha$ by Hameed & Devereux (1999). The $H\alpha$ image shows emission from H_{II} regions in the spiral arms, weak diffuse emission in the bulge, a nuclear source and a ring-like structure surrounding the nucleus. The GMRT radio continuum morphology (Fig. 4) shows a kpc-scale linear jet like structure and a bright radio source coincident with the $H\alpha$ nuclear source. No $H\alpha$ emission is seen corresponding to the extended radio emission. Therefore, it is unlikely that the extended radio emission is due to the star formation activities, and could be due to a radio jet. The radio morphology, the presence of a nuclear point source in radio and in $H\alpha$ and a significant radio excess indicate an AGN in this galaxy. The apparent alignment of the linear radio structure with the major axis of NGC 1371 is likely to be due to a projection effect although in reality the jet may be at a considerable angle from the plane of the galaxy.

6. The normal star-forming galaxies

The majority (70%) of galaxies in the Eridanus group have their star formation rates (SFR) below that of the Milky Way, viz., $\sim 1 M_\odot \text{ yr}^{-1}$. The SFR can be inferred from the FIR luminosity or from the radio luminosity. The SFR (α) can be estimated from the FIR flux densities using the relation (Kennicutt 1998):

$$\alpha = 1.7 \times 10^{-10} f (1 + 0.387 S_{100\mu\text{m}} / S_{60\mu\text{m}}) L_{60\mu\text{m}}. \quad (5)$$

The value of f is quite uncertain and varies in the range 1.2–2 for star-burst galaxies (Sanders *et al.* 1991).

The value of α can be estimated using the 20 cm radio spectral luminosity (Yun *et al.* 2001):

$$\alpha \sim 6 \times 10^{-22} \frac{L_{20\text{ cm}}}{\text{W Hz}^{-1}}. \quad (6)$$

The radio continuum images of some of the normal star forming galaxies in the Eridanus group are shown in Fig. 5. The radio emission is diffuse in these galaxies. It can be seen that the radio continuum emission in NGC 1309 and NGC 1385 follows the spiral arms. This is a common trend observed in spiral galaxies where it is believed that the spiral density wave compresses the magnetic field in the spiral arms thereby increasing the magnetic field strength and hence the synchrotron emission.

The most FIR luminous galaxies in the group are NGC 1482 ($L_{\text{FIR}} \sim 4.3 \times 10^{10} L_{\odot}$) and NGC 1385 ($L_{\text{FIR}} \sim 2.5 \times 10^{10} L_{\odot}$). Although these galaxies are ‘star-bursts’ (Table 3), they are called ‘normal’ since they follow the radio–FIR correlation. Both of these galaxies have signatures of recent tidal interactions in their optical images. NGC 1385 (SBcd) has a faint stellar envelop to the south and disturbed spiral arms. NGC 1482 (SA0) has faint stellar streamers surrounding it. The H_I tidal tails were detected in NGC 1482 (Omar & Dwarakanath 2005b). The large amounts ($M_{\text{H}_2} \sim 10^9 M_{\odot}$) of molecular hydrogen in the centres of these two galaxies have been inferred from the CO observations (Sanders *et al.* 1991; Elfhag *et al.* 1996). Strong stellar winds indicative of a star-burst were detected in H $_{\alpha}$ (Hameed & Devereux 1999; Veilleux & Rupke 2002).

7. The FIR-excess (radio-deficient) galaxies

Galaxies having significant excess FIR emission are not common. Yun *et al.* (2001) noticed 9 galaxies with q value greater than 3 in a sample of nearly 1800 galaxies. A significantly higher q value implies that the galaxy has excess FIR emission than expected from the radio luminosity. Galaxies with $L_{60\mu} \sim 10^8 L_{\odot}$ are likely to have excess FIR emission due to the heating of dust from the old stellar population (Condon *et al.* 1991). These galaxies are not considered here. The nine FIR excess galaxies of Yun *et al.* (2001) have $L_{60\mu} > 10^9 L_{\odot}$. The FIR excess in these galaxies was diagnosed recently by Roussel *et al.* (2003) who showed that these galaxies are most likely observed within a few Myr of the onset of an intense star formation episode after being quiescent for at least 100 Myr. According to them, this star-burst, while heating the dust, has not produced cosmic rays yet. Therefore, these galaxies are radio-deficient. Such galaxies must be rare to be observed at a given epoch since the typical ages of massive stars (~ 1 Myr) are only a small fraction of the ages of the star-bursts (~ 100 Myr) in galaxies. The actual number of such synchrotron deficient galaxies will depend further on the time intervals of successive star-burst events in a galaxy.

The Eridanus group has two such galaxies, viz., NGC 1377 and IC 1953. Both these galaxies were studied by Roussel *et al.* (2003). The q values are ~ 3.4 for NGC 1377 and ~ 2.9 for IC 1953. These two galaxies are marked in Fig. 3. A brief description of these two galaxies is given below.

7.1 NGC 1377 (S0)

NGC 1377 is FIR luminous ($L_{\text{FIR}} = 8 \times 10^9 L_{\odot}$) with $S_{60\mu}/S_{100\mu} \sim 1.3$ indicating the presence of warm dust in the galaxy. The warm IRAS galaxies are believed to be

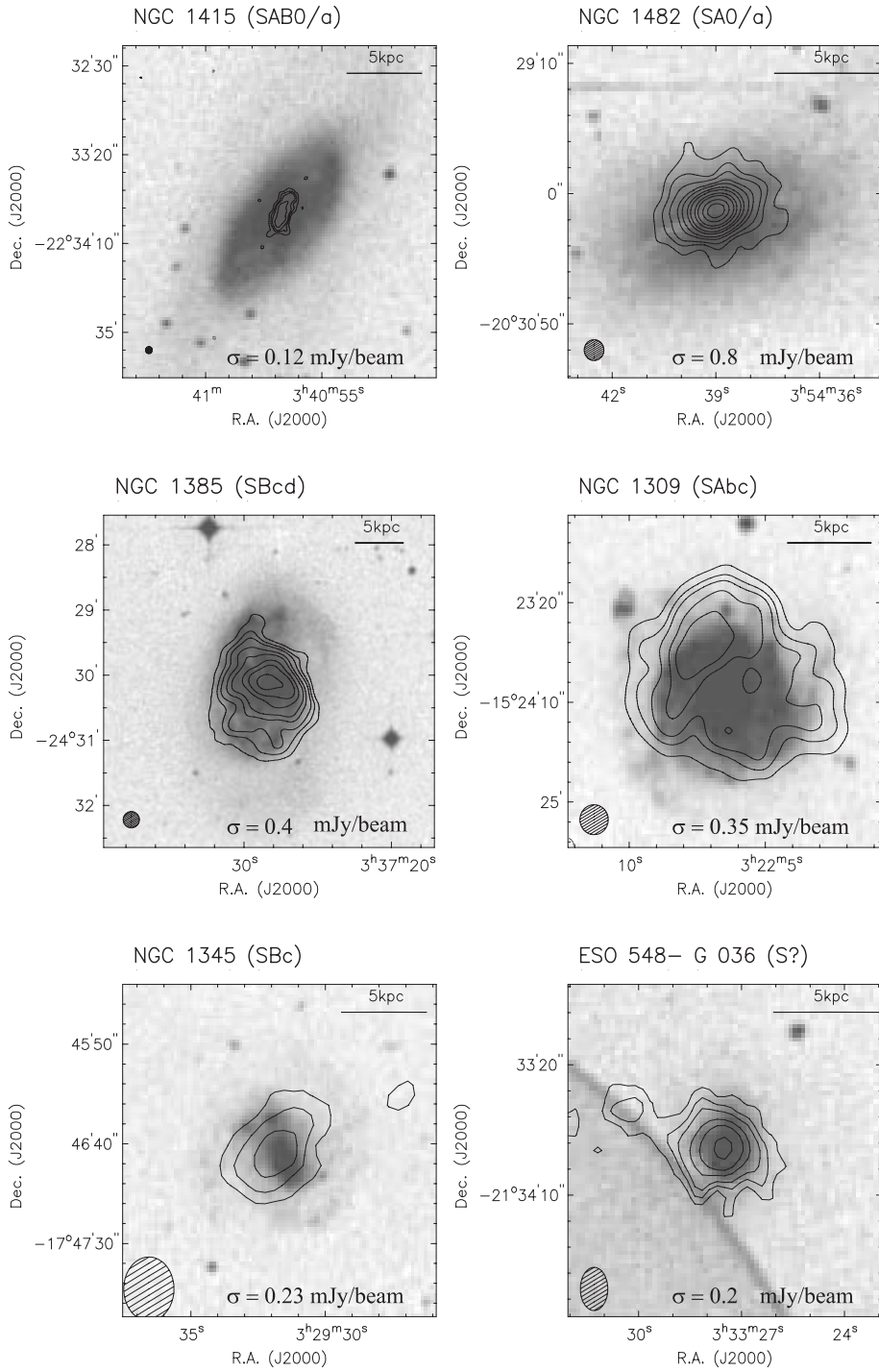


Figure 5. GMRT radio continuum images (in contours) of some of the normal star forming galaxies in the Eridanus group overlaid upon their respective optical images from the DSS (see Table 3). The contour levels are similar to Fig. 4. The rms values of the respective images are indicated in each image.

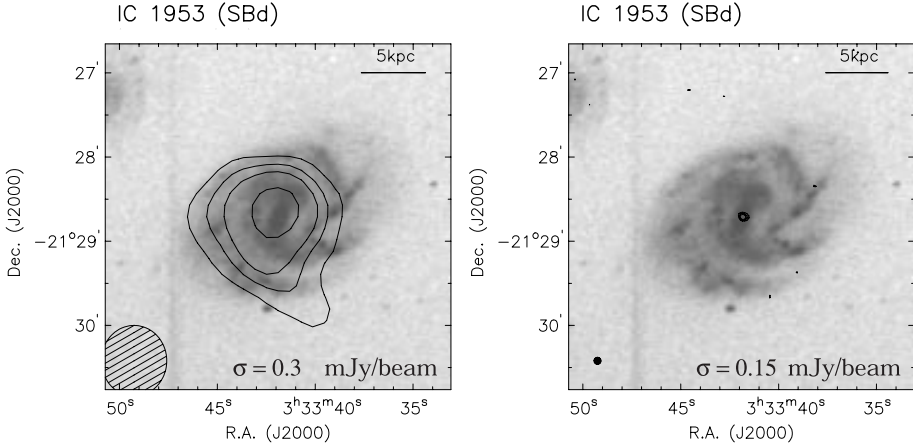


Figure 6. The GMRT radio continuum morphologies (in contours) of IC 1953 at resolutions of $50'' \times 45''$ (left) and $5'' \times 5''$ (right) overlaid on the optical image from the DSS. The contours are at 0.9, 1.4, 1.8 and $2.7 \text{ mJy beam}^{-1}$ for the lower resolution image and 0.5, 0.7, and $0.9 \text{ mJy beam}^{-1}$ for the higher resolution image.

due to nuclear star-burst probably triggered by recent interactions or mergers (Heisler & Vader 1995). The radio continuum emission from this galaxy is not detected in the GMRT image down to a 5σ sensitivity limit of 1 mJy . NGC 1377 is, therefore, radio deficient by at least 40 times.

7.2 IC 1953 (SBd)

IC 1953 has grand spiral arms and a prominent bar. The radio continuum morphology consists of diffuse emission from the disk and a nuclear point source (Fig. 6). Using high resolution (kpc-scale) mid-infrared images from ISO (Infrared Space Observatory) observations and high resolution radio continuum images from the VLA observations, Roussel *et al.* (2003) showed that while the disk radio continuum emission in IC 1953 is consistent with the disk infrared emission, the nuclear radio emission is not. According to them, it is suggestive of a recent nuclear star-burst.

8. Conclusions

- The upper end of the radio luminosity distribution of the Eridanus galaxies ($L_{20 \text{ cm}} \sim 10^{20} \text{ W Hz}^{-1}$) is consistent with that of field galaxies, other groups, and late-type galaxies in clusters of galaxies.
- The Eridanus group lacks powerful AGNs ($L_{20 \text{ cm}} > 10^{23} \text{ W Hz}^{-1}$) commonly found in rich clusters.
- Most of the Eridanus galaxies follow the well-known radio–FIR correlation.
- Majority ($\sim 70\%$) of the Eridanus galaxies have their SFR below that of the Milky Way.
- Two previously unknown low luminosity AGNs ($L_{20 \text{ cm}} < 10^{22} \text{ W Hz}^{-1}$) have been detected in the group.

- The most FIR and radio luminous galaxies in the Eridanus group have their optical and H_I morphologies suggestive of recent tidal interactions.
- The Eridanus group has two FIR luminous ($L_{\text{FIR}} \sim 10^{10} L_{\odot}$) but radio-deficient galaxies. It is believed that these galaxies are most likely observed within a few Myr of the onset of an intense star formation episode after being quiescent for at least 100 Myr.

Acknowledgements

We thank the staff of the GMRT who made these observations possible. The GMRT is operated by the National Centre for Radio Astrophysics of the Tata Institute of Fundamental Research. This research has made use of the NASA/IPAC Infrared archived data and Northern VLA Sky Survey data. This research has been benefited by NASA's Astrophysics Data System (ADS) and Extra-galactic Database (NED) services.

Appendix A: The FIR and radio continuum properties of the Eridanus group of galaxies

Table 2. FIR and radio properties of Eridanus galaxies.

Name	H.T.	$S_{60\mu}$ (Jy)	$S_{100\mu}$ (Jy)	$\log \frac{L_{\text{FIR}}}{L_{\odot}}$	$S_{20\text{ cm}}$ (mJy)	$\log \frac{L_{20\text{ cm}}}{\text{WHz}^{-1}}$	SFR $M_{\odot}/\text{yr}$
NGC 1300	Sbbc	2.75	10.3	9.77	52.1	21.52	1.96
UGCA 068	SABm	0.21	0.56	8.57	–	< 20	< 0.06
NGC 1325	SAbc	0.63	3.21	9.21	–	< 20	< 0.06
NGC 1325A	SBbc	0.24	0.85	8.70	–	< 20	< 0.06
NGC 1332	S0	0.50	1.78	9.02	4.0	20.40	0.15
NGC 1345	SBc	0.78	1.59	9.06	3.9	20.40	0.15
NGC 1347	?	0.30	0.99	8.78	–	< 20	< 0.06
ESO 548-G 028	SB0	0.29	0.61	8.66	–	< 20	< 0.06
ESO 548-G 029	SB?	0.24	0.73	8.66	–	< 20	< 0.06
NGC 1353	SAbc	2.42	8.79	9.70	5.5	20.54	0.21
IC 1952	SBbc	0.79	3.19	9.25	7.2	20.66	0.27
IC 1953	SBd	8.47	11.28	10.05	13.0	20.92	0.49
NGC 1359	SBm	2.13	4.28	9.52	32.2	21.31	1.21
ESO 548-G 047	SB0	0.39	1.00	8.83	–	< 20	< 0.06
NGC 1370	E+	1.19	2.17	9.25	3.5	20.35	0.13
NGC 1377	S0	7.25	5.74	9.92	–	< 20	< 0.06
NGC 1385	SBcd	15.87	33.78	10.40	190	22.08	7.13
NGC 1395	E2	0.05	0.34	8.20	–	< 20	< 0.06
NGC 1400	S0	0.72	2.50	9.17	1.7	20.03	0.06

Table 2. (Continued)

Name	H.T.	$S_{60\mu}$ (Jy)	$S_{100\mu}$ (Jy)	$\log \frac{L_{\text{FIR}}}{L_{\odot}}$	$S_{20\text{ cm}}$ (mJy)	$\log \frac{L_{20\text{ cm}}}{\text{WHz}^{-1}}$	SFR $M_{\odot}/\text{yr}$
NGC 1407	E0	0.14	0.48	8.45	93.3	21.77	3.50
NGC 1415	SAB0/a	5.27	12.73	9.95	25.8	21.22	0.97
ESO 482-G 035	SBab	0.26	1.00	8.75	–	< 20	< 0.06
NGC 1422	SBab	0.39	1.09	8.85	–	< 20	< 0.06
NGC 1439	E1	–	0.34	–	–	–	< 20
NGC 1440	SB0	–	1.29	–	–	–	< 20
NGC 1438	SB0/a	0.23	0.79	8.67	–	< 20	< 0.06
NGC 1481	SA0	0.36	–	8.50	3.7	20.37	0.14
ESO 548-G 036	S?	1.94	–	9.23	7.4	20.67	0.28
ESO 548-G 043	Sa	0.35	–	8.49	–	< 20	< 0.06
UGCA 073	E+	0.3	–	8.42	–	< 20	< 0.06
NGC 1145	Sc	0.98	3.50	9.31	4.5	20.46	0.17
NGC 1163	SBbc	0.64	1.71	9.06	2.3	20.17	0.09
MCG-03-08-057	SAd	0.57	2.39	9.12	–	< 20	< 0.06
NGC 1187	SBc	10.54	22.41	10.22	69.6	21.65	2.61
NGC 1179	SBd	0.41	2.21	9.04	–	< 20	< 0.06
NGC 1231	Sc	0.23	0.58	8.60	–	< 20	< 0.06
MRK 1069	?	1.28	1.86	9.24	9.5	20.78	0.36
NGC 1232	SABc	3.47	21.7	10.02	75.5	21.68	2.83
IC 1898	SBc	0.78	2.77	9.21	5.5	20.54	0.21
NGC 1255	SABbc	2.99	11.36	9.81	38.0	21.38	1.43
NGC 1258	SABcd	0.28	1.09	8.79	–	< 20	< 0.06
NGC 1292	SAc	1.36	4.38	9.43	4.4	20.45	0.17
UGCA 064	SBcd	0.19	0.91	8.68	–	< 20	< 0.06
NGC 1302	SABa	0.26	1.74	8.91	–	< 20	< 0.06
NGC 1306	Sb	2.63	4.73	9.59	15.2	20.99	0.57
IRAS 03191-2456	S0	0.53	0.78	8.86	–	< 20	< 0.06
NGC 1309	SAbc	5.22	14.26	9.97	75.2	21.68	2.82
UGCA 071	Sd	0.28	0.81	8.71	2.76	20.24	0.10
NGC 1338	SABb	1.36	4.95	9.46	12.3	20.89	0.46
ESO 481-G 029	S0	0.48	1.16	8.91	–	< 20	< 0.06
ESO 418-G 008	SBd	0.48	1.22	8.92	3.9	20.40	0.15
NGC 1357	SAab	0.93	4.67	9.38	4.4	20.45	0.17
NGC 1371	SABa	0.20	1.36	8.80	15.8	21.00	0.59
NGC 1398	SBab	1.14	8.96	9.60	29.0	21.27	1.09
NGC 1425	SAb	1.08	5.89	9.47	26.6	21.23	1.00

Table 2. (Continued)

Name	H.T.	$S_{60\mu}$ (Jy)	$S_{100\mu}$ (Jy)	$\log \frac{L_{\text{FIR}}}{L_{\odot}}$	$S_{20\text{ cm}}$ (mJy)	$\log \frac{L_{20\text{ cm}}}{\text{WHz}^{-1}}$	SFR $\text{M}_{\odot}/\text{yr}$
NGC 1421	SABbc	8.48	21.32	10.16	114.5	21.86	4.30
MCG-02-10-009	Sc	0.53	2.13	9.07	–	< 20	< 0.06
MCG-03-10-042	SABbc	0.86	3.40	9.28	9.1	20.76	0.34
J034559-1231	?	0.63	1.68	9.05		< 20	< 0.06
MCG -03-10-045	IBm	4.77	7.88	9.83	22.6	21.16	0.85
NGC 1461	SA0	0.08	0.31	8.24	–	< 20	< 0.06
UGCA 085	Sc	0.24	0.93	8.72	–	< 20	< 0.06
NGC 1464	?	2.61	4.89	9.59	21.1	21.13	0.79
NGC 1482	SA0	31.95	45.32	10.63	243	22.19	9.12
IC 2007	SBc	1.28	2.77	9.31	5.6	20.55	0.21
NGC 1518	SBdm	2.16	6.55	9.61	32.3	21.31	1.21
NGC 1519	SBb	0.91	2.48	9.21	–	< 20	< 0.06
ESO 483-G 013	SA0	0.39	1.08	8.85	2.0	20.11	0.08
J041343-1729	?	0.98	1.77	9.16		< 20	< 0.06
ESO 550-G 008	E?	0.17	–	8.17	–	< 20	< 0.06
UGCA 061	SBm	0.19	–	8.22	–	< 20	< 0.06
IRAS 03007-1531	?	0.36	–	8.50	–	< 20	< 0.06

Column 1 – Name of the galaxy; **Column 2** – Hubble type; **Column 3** – 60μ flux density from IRAS; **Column 4** – 100μ flux density from IRAS; **Column 5** – Total FIR luminosity using equation 2; **Column 6** – 20 cm flux density from NVSS; **Column 7** – 20 cm spectral luminosity using equation 3; **Column 8** – Star formation rate using equation 6.

Table 3. Galaxies studied with the GMRT.

Name	Type	q	$S_{60\mu}/S_{100\mu}$	Comments
NGC 1407	E	0.45	0.29	AGN, radio-excess
NGC 1371	Sa	1.8	0.15	AGN, radio-excess
NGC 1415	SB0/a	2.5	0.41	Normal
NGC 1482	S0/a	2.3	0.70	Star-burst
NGC 1385	SBcd	2.1	0.47	Star-burst
NGC 1377	S0	> 3.9	1.26	Radio-deficient
IC 1953	SBcd	3.1	0.75	Radio-deficient
NGC 1309	Sbc	1.95	0.37	Normal
NGC 1345	Sbc	2.43	0.49	Normal
ESO 548-G 036	S?	2.08	–	Normal

References

- Condon, J. J. 1989, *Astrophys. J.*, **338**, 13.
- Condon, J. J., Anderson, M. L., Helou, G. 1991, *Astrophys. J.*, **376**, 95.
- Condon, J. J. 1992, *Ann. Rev. Astron. Astrophys.*, **30**, 575.
- Condon, J. J., Cotton, W. D., Greisen, E. W. *et al.* 1998, *Astron. J.*, **115**, 1693.
- Elfagh, T., Booth, R. S., Hoeglund, B. *et al.* 1996, *Astron. Astrophys. Supl.*, **115**, 439.
- Gavazzi, G., Boselli, A. 1999, *Astron. Astrophys.*, **343**, 93.
- Helou, G., Khan, I. R., Malek, L., Boehmer, L. 1988, *Astrophys. J. Supl.*, **68**, 151.
- Hameed, S., Devereux, N. 1999, *Astron. J.*, **118**, 730.
- Harwit, M., Pacini, F. 1975, *Astrophys. J. Lett.*, **200**, 127.
- Heisler, C. A., Vader, J. P. 1995, *Astron. J.*, **110**, 87.
- Kennicutt, R. C. 1998, *Ann. Rev. Astron. Astrophys.*, **36**, 189.
- Menon, T. K. 1995, *Mon. Not. R. Astron. Soc.*, **274**, 845.
- Miller, N., Owen, F. N. 2001, *Astron. J.*, **121**, 1903.
- Neugebauer, G. *et al.* 1984, *Astrophys. J. Lett.*, **278**, 1.
- Omar, A. 2004, Ph.D. thesis, Jawaharlal Nehru University, Delhi.
- Omar, A., Dwarakanath, K. S. 2005a, *J. Astrophys. Astr.*, this volume 1.
- Omar, A., Dwarakanath, K. S. 2005b, *J. Astrophys. Astr.*, this volume 71.
- Reddy, N. A., Yun, M. S. 2004, *Astrophys. J.*, **600**, 695.
- Roussel, H., Helou, G., Beck, R. *et al.* 2003, *Astrophys. J.*, **593**, 733.
- Sanders, D. B., Scoville, N. Z., Soifer, B. T. 1991, *Astrophys. J.*, **370**, 158.
- Scoddeggio, M., Gavazzi, G. 1993, *Astrophys. J.*, **409**, 110.
- van der Kruit, P. C. 1973, *Astron. Astrophys.*, **29** 263.
- Veilleux, S., Rupke, D. S. 2002, *Astrophys. J. Lett.*, **565**, 63.
- Verheijen, M. A. W., Sancisi, R. 2001, *Astron. Astrophys.*, **370**, 765.
- Willmer, C. N. A., Focardi, P., da Costa, L. N., Pellegrini, P. S. 1989, *Astron. J.*, **98**, 1531.
- Yun, M. S., Reddy, N. A., Condon, J. J. 2001, *Astrophys. J.*, **554**, 803.

Parametric Identification of Solar Series based on an Adaptive Parallel Methodology

Juan A. Gómez Pulido¹, Miguel A. Vega Rodríguez & Juan M. Sánchez Pérez

Department of Computer Science, University of Extremadura, Campus Universitario s/n. 10071 Cáceres, Spain.

¹*e-mail: jangomez@unex.es*

Received 2004 September 4; accepted 2005 March 21

Abstract. In this work we present an adaptive parallel methodology to optimize the identification of time series through parametric models, applying it to the case of sunspot series. We employ high precision computation of system identification algorithms, and use recursive least squares processing and ARMAX (Autoregressive Moving Average Extensive) parametric modelling. This methodology could be very useful when the high precision mathematical modelling of dynamic complex systems is required. After explaining the proposed heuristics and the tuning of its parameters, we show the results we have found for several solar series using different implementations. Thus, we demonstrate how the result precision improves.

Key words. Sunspot—time series—system identification—parametric modelling—optimization, parallelism—adaptation.

1. Introduction

In many science and engineering fields, it is necessary to build mathematical models for studying the behaviour of phenomena and systems whose mathematical description is not available “*a priori*”. One interesting type of these systems is the Time Series (TS). Time series are used to describe behaviours in many fields: astrophysics, meteorology, economy, etc. When dealing with TS there is only one signal available under observation; its physical structure is not known. This led us to employ the planning System Identification (SI) techniques (Söderström *et al.* 1989) in order to obtain the TS model. The model precision depends on the assigned values to certain parameters.

In this paper we have focused the effort of analysis in a kind of solar time series: the sunspot series. Then, we propose a parallel and adaptive heuristics to adjust the system identification main parameters with the aim of improving the precision of the parametric model of the sunspot series.

1.1 Time series modelling

We consider TS as a sampled signal with period T that is modelled with an ARMAX (Ljung 1999) parametric polynomial description (see equation 1).

$$y(k) + a_1 y(k_1) + \dots + a_{na} y(k_{na}) = 0 \quad \text{where } na = \text{dimension.} \quad (1)$$

Basically the identification consists in determining the ARMAX model parameters a_i (θ in matricial notation) from measured samples $y(k_i)$ ($\varphi(k)$ in matricial notation). Then it is possible to compute the estimated signal $y_e(k)$ (equation 2) and compare it with the real signal $y(k)$, computing the generated error (equation 3)

$$y_e(k) = [-a_1 y(k-1) - \dots - a_{na} y(k-na)] = \varphi^T(k) \theta, \quad (2)$$

$$\text{error}(k) = y(k) - y_e(k). \quad (3)$$

The recursive estimation updates a_i in each time step k , thus modelling the system. The more sampled data processed, the more precision for the model, because it has more information about the system behaviour history. We consider SI performed by the well-known Recursive Least Squares (RLS) with forgetting factor (λ) algorithm (Ljung 1999). From the initial conditions, we start building $\varphi^T(k)$, and then RLS follows as it is shown in Fig. 1.

This algorithm is specified by the constant λ (forgetting factor), the initial values and the observed samples $\{y(k)\}$. There is no fixed value for λ , even if it is used as a value between 0.97 and 0.995 (Ljung 1991). The cost function F (see equation 4) is defined as the value to minimize in order to obtain the best precision.

$$F(\lambda) = \sum_{k=k_0}^{k=k_0+SN-1} |y_e(k) - y(k)| \quad (SN \text{ is the sample number}). \quad (4)$$

The recursive identification is very useful when it is a matter of predicting the following behaviour of the time series from the data observed up to the moment. For many purposes, it is necessary to make this prediction, and for predicting it is necessary to obtain information about the system. This information, acquired by means of the SI, consists in elaborating a mathematical parametric model for covering the system behaviour.

SI allows finding, in sample time, a mathematical model ($\theta(k)$) from which it is possible to predict the next behaviours. As identification advances in time, the

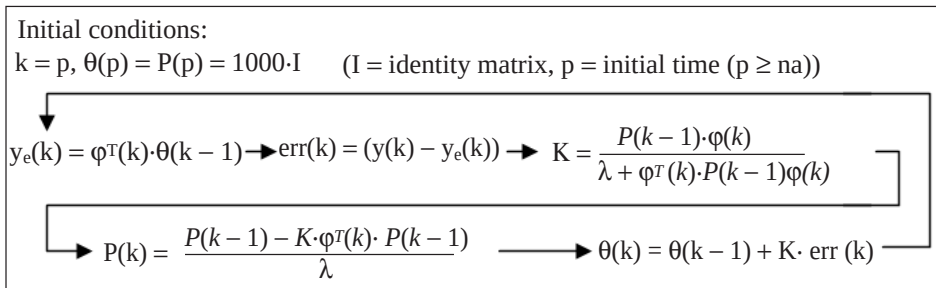


Figure 1. RLS equations.

predictions improve using more precise models. For example, we can compute in sample time the system model and then, with this model simulate the system future behaviour, forwarding real situations (Fig. 2).

For validating the SI evolutionary parallel architecture that we present, we use a time series set, corresponding to sunspot series obtained from measured observations (ROB 2004) (NOAA 2004). We have used 13 time series (Fig. 3) showing daily sunspots: 10 series (ss_00, ss_10, ss_20, ss_30, ss_40, ss_50, ss_60, ss_70, ss_80 and ss_90) each one corresponding to the sunspot measurements during ten years (for example, ss_20 compiles the sunspots from 1/1/1920 to 31/12/1929); 2 series (ss_00_40 and ss_50_90) each one covering 50 years, and 1 series (ss_00_90) covering all measurements of the 20th century.

2. The optimization problem

When SI techniques are used, the model is generated “*a posteriori*” by means of the measured data. However, we are interested in the system behaviour prediction in running time, that is, while the system is working and its data are being observed. So, it would be interesting to generate models in running time in such a way that a processor may simulate the system next behaviour.

At the same time, our first effort is to obtain a high model precision (minimal F). SI precision is due to several causes, mainly to the forgetting factor λ (Fig. 4). Frequently this value is critical for model precision. Other sources can also have lesser degree of influence (dimensions, initial values, the system...), but they are considered as problem definitions, not parameters to be optimized.

On the other hand, the precision problem may appear when a system model is generated in sample time: If the system response changes quickly, then the sample frequency must be high for avoiding the key data loss in the system behaviour description. If the system is complex and its simulation from the model to be found must be very trustworthy, then the required precision must be very high and this implies a great computational cost. Sometimes the hardware resources do not allow the computational cost in the model generation and processing to be lower than the sample period. We find a trade-off between a high sample frequency and a high precision in the algorithm computation. Adjusting the parameter λ for improving the precision can be conveniently done by means of techniques based on adaptive heuristics. So, the final goal is to design an architecture for real time SI suited for processes in which a high precision and a low sample time are required.

The method we suggest, and explain in the next section, can be used in any area. However, for the case of solar time series, the interval of sampling is not important, the real goal being to obtain the maximum accuracy of the estimation.

3. PARLS: the proposed heuristics

In order to find the optimum value of λ , we propose a parallel algorithm that is partially inspired on the concept of artificial evolution (Goldberg 1989) (Rechenberg 1973) and also in simulated annealing mechanism (Kirkpatrick *et al.* 1983). In our algorithm, named PARLS (Parallel Adaptive Recursive Least Squares), the optimization parameter λ is evolved for predicting new situations during the successive phases of the

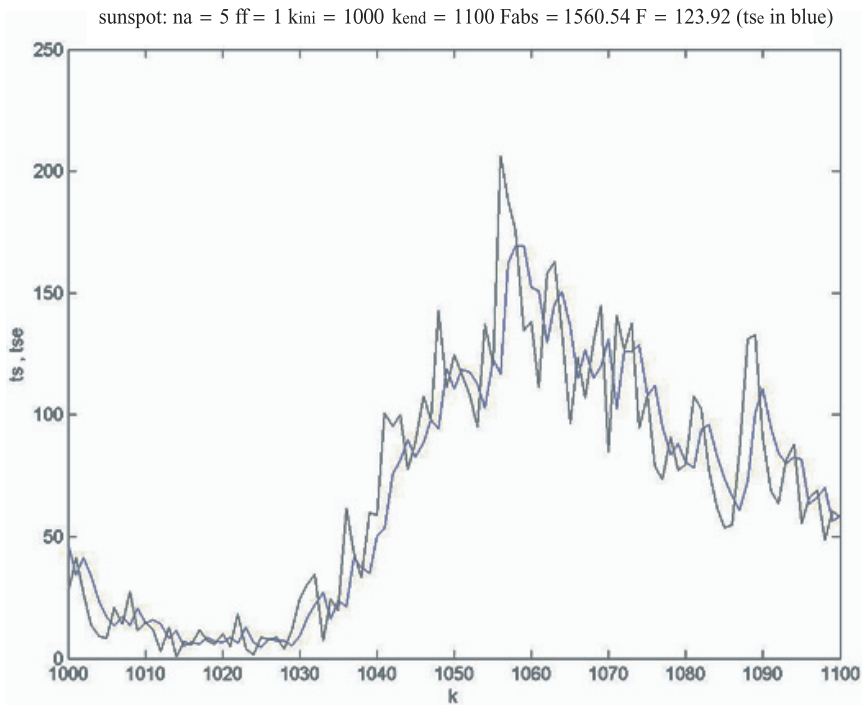
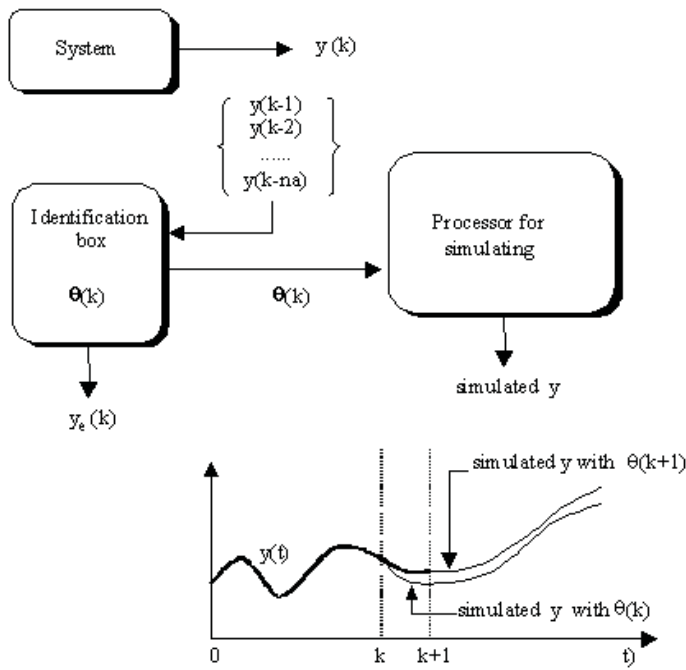


Figure 2. Recursive SI allows us to predict and simulate future system behaviour. On the right side, details of a sunspot time series used and its simulation based on the model identified.

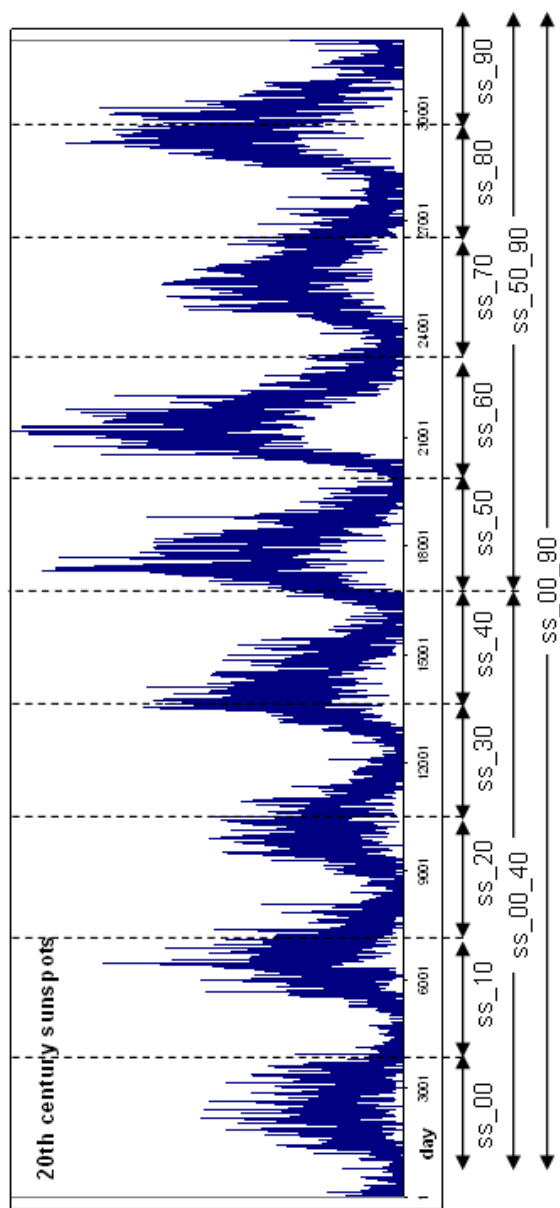


Figure 3. The sunspot time series used in this work.

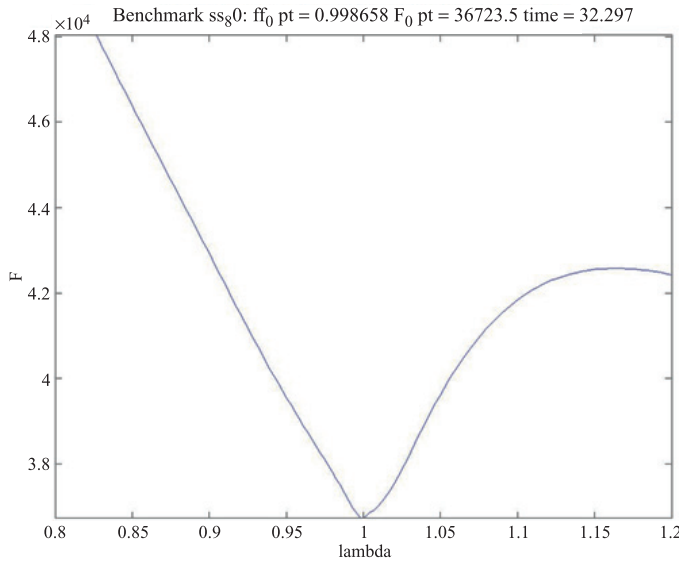


Figure 4. Cost function evaluated for several λ values using RLS identification for ss_80 benchmark when $na = 5$. We can see how the best performance is reached when $\lambda = 0.998658$.

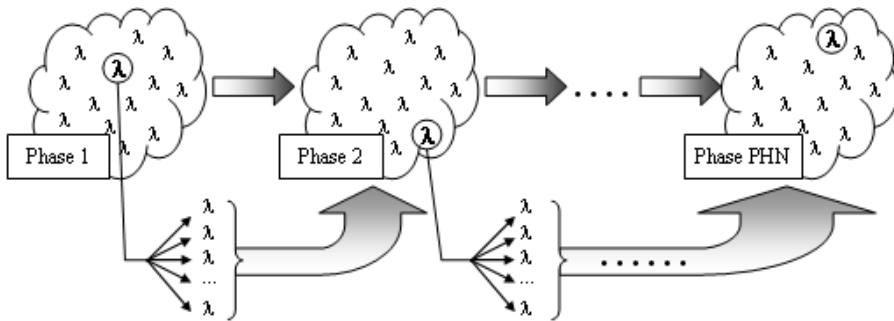


Figure 5. Evolution mechanism. In each phase, a set of λ values performs RLS identification during a certain number of samples. Then, the λ value whose corresponding F is the minimum of all computed F is the optimum, and from it a new set of λ values is generated. They are used in the next PARLS phase to perform new identification during the following PHS samples.

process (iterations of the algorithm). In other words, λ evolves at the same time that improves the cost function performance (Fig. 5).

PARLS considers a λ value as a state. Starting on an initial λ value (λ_c) and an initial R value (the interval of generation where λ_c is in the middle), a set of λ values is generated covering the entire interval R uniformly. The λ values generated are equal to the number of parallel processing units (PUN). Each phase of PARLS process is an identification loop that considers a given number of sample times (PHS) and the corresponding λ value. In this work, we use the nomenclature shown in Table 1.

In each phase, R is reduced dividing itself by the RED factor (the interval limits are moved so that the center of the interval corresponds with the optimal λ value found in the previous phase), in such a way that the generated set of λ will be more and

Table 1. PARLS nomenclature.

R	Generation interval
λ_c	λ central in R
PHS	Phase samples
PHN	Number of phases
PUN	Number of parallel processing units
TSN	Total number of samples
RED	Reduction factor of R

more near to the previous optimum found. The new set of generated λ values always covers uniformly the new interval R . In each processing unit, during each phase, the cost function F is computed (F is defined as the accumulated error of the samples that constitute each phase). From equation (4), we have:

$$F(\lambda_{\text{PUN}}) = \sum_{k=k_0}^{k=k_0+PHS-1} |y_e(k) - y(k)|. \quad (5)$$

At the end of each phase, the best λ is chosen. This is the corresponding value to the lower F . From this λ , new values are generated in a more reduced (new R) interval (Figs. 6, 7). The goal is that the identifications performed by the processing units will converge to optimum λ parameters when a given stop criteria is achieved. So the identification will be of high precision.

Therefore, PARLS could be considered as population-based metaheuristics rather than parallel metaheuristics, because each processing unit is able to operate in isolation, as well as the tackled problem itself as only a single real-valued parameter (λ) is optimized.

4. Experimental results

We consider several criteria for evaluating PARLS performances. All these criteria have been fully checked and tested (Gómez *et al.* 2001) in order to get a set of better values for parameters and strategies. For example, we have studied strategies as the optimum λ criteria (the λ value that produces a minimum F), the stop criteria (indicating when a processing unit must stop the work), the model generation criteria (how to consider the initial model in the next phase), the optimum F definition (to consider the optimum F as the lowest in all phases or the lowest computed in the present phase), etc.

A question of interest is about the optimal model size. There is a greater computational cost when na is higher. Therefore, the na value should be selected in relation to the required computational cost, and this dimension is considered as part of the definition of the problem. We have computed several problems with different values for na , using SISO (Single Input Single Output) systems instead of time series and with processing units implemented with neural network (Gomez *et al.* 2002), and our conclusion is that an accurate tuning for general purposes could be from $na = 5$, although this value can be increased a lot for systems with wide sampling periods. Because a bigger na increases the model precision, for the sunspot series an adequate value of

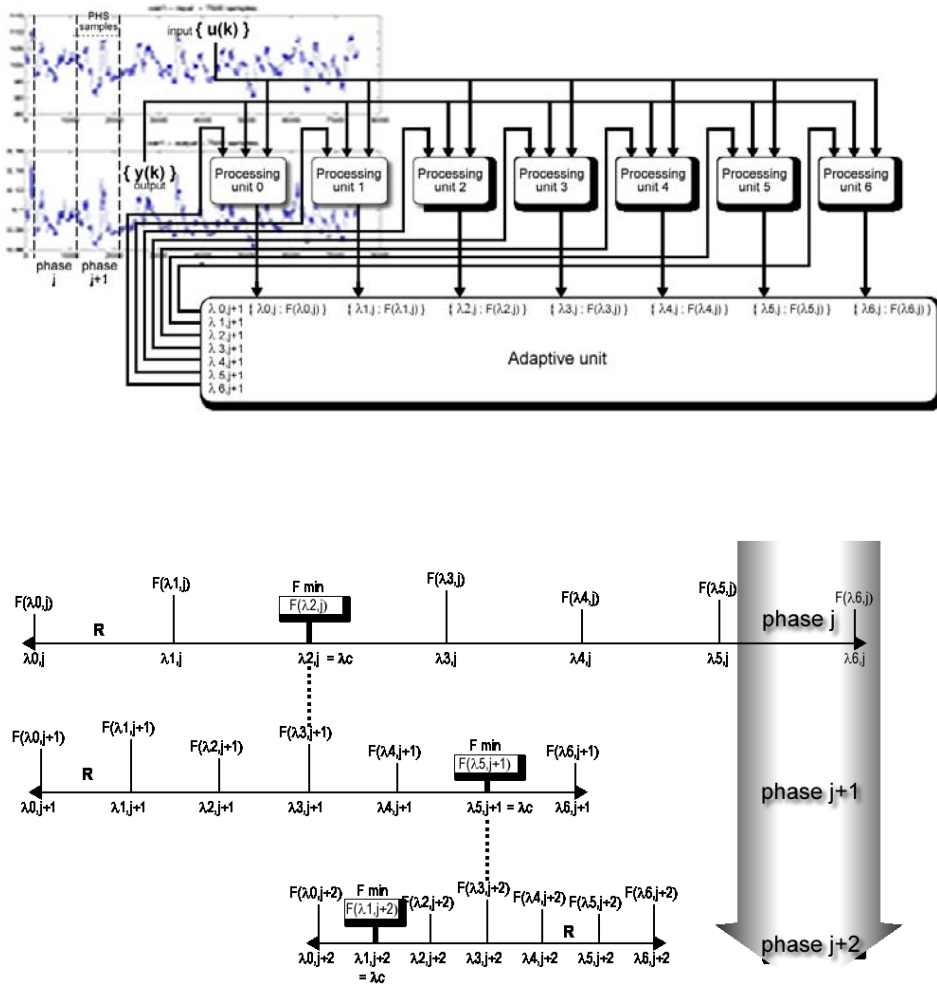


Figure 6. The SI uses different λ values in each phase performed by each processing unit. All the λ values in the same phase running in the processing units are generated in the R interval from the previous phase optimum λ found, corresponding with the smallest computed F .

na can be 20, since the computational cost is not big when they are series with a high interval of sampling (one day).

Another important question is how to establish the initial range of λ values in PARLS search. We have performed experiments in order to determine the approximated optimal λ for several series using a lot of RLS computations. In Fig. 8, some of these experiments are shown. We can see that there is a distinct optimal λ for each series, but in all the cases there is a smooth V-curve that is very useful to the initial PARLS search. We have thus selected as initial searching parameters tuned values $\lambda_c = 1$ and $R = 0.05$. In view of the results, it could be said that the optimum λ value is 1, but it is dangerous to state this if we have not reduced the range more. Thus, reducing the interval of search, we will be able to find a near, but different optimum value to 1, as we can see in Fig. 9.

```

 $\lambda c$  /*  $\lambda$  value in the middle of the set of  $\lambda$  values in the interval */
R /* Length of that interval */
do {
  Generating PUN  $\lambda$ 's from  $\lambda c$ , covering all R interval
  do { /* phase = Identification of PHS samples */
    Computing accumulated error (F) in
    each PU during all samples of the phase
  }
  while (processing PHS samples);
  F( $\lambda$ ) = Accumulated error for each PU
  Determining  $\lambda$  optimum of all PUs for which F( $\lambda$ ) is minimum
   $\lambda c = \lambda_{opt}$ ;
  R = R/RED; /* R is reduced in a RED factor */
}
while (!stop criterion)

```

Figure 7. PARLS pseudo-code.

PARLS offers a great variability for its parameters. According to the results we have obtained, we can conclude that there are no common policies for tuning the parameters in such a way that the best results will always be found. But results indicate that there is a set of values for which the results are good. We can thus establish fixed values for PARLS parameters (Table 2) in order to define a unique algorithm applicable to any series. This has the advantage of a quick application for a series without the previous task of tuning parameters.

5. Conclusions

In Table 3, we show the comparison of found results between RLS search and PARLS heuristics for several sunspot series. In all the cases the same tuned parameters have been used ($na = 20$, $PUN = 11$, $\lambda c = 1$, $R = 0.05$, $RED = 2$, $PHN = 4$). In (a) the results of 11 RLS identifications with their corresponding 11 equidistant in R values of λ are shown, and in (b) the PARLS results are displayed too. The computational effort of 11 RLS identifications is almost equal to PARLS cost with 11 processing units, so both results can be compared in order to establish the conclusions. With these tuned parameters, PARLS always finds better results than RLS. This fact contrasts with the results obtained for series of other areas (Gomez *et al.* 2002), for which PARLS finds better results in most of them, but for a few series the difference of F oscillates between 2% and 5% RLS better than PARLS. Therefore, if for the series studied in Gomez *et al.* (2002) we said that PARLS improves or holds the results found with RLS with the same computational effort, now, for the sunspot solar series, we can say PARLS always gives us better results.

As starting point, we can say that the parallel adaptive heuristic PARLS offers a good performance, and this encourages us to follow this research. Also, nowadays, we are trying to accelerate the computation of some parts of PARLS using reconfigurable hardware to obtain a synthesis on specialized coprocessors capable of improving the global efficiency (Gomez *et al.* 2002) for series of other areas where the sampling interval is very short.

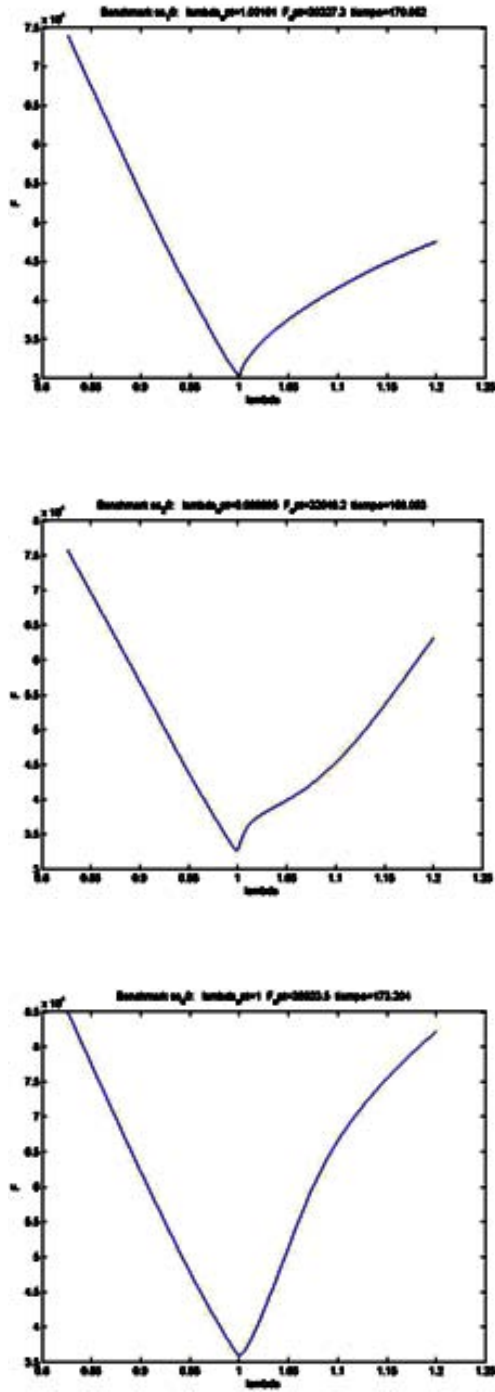
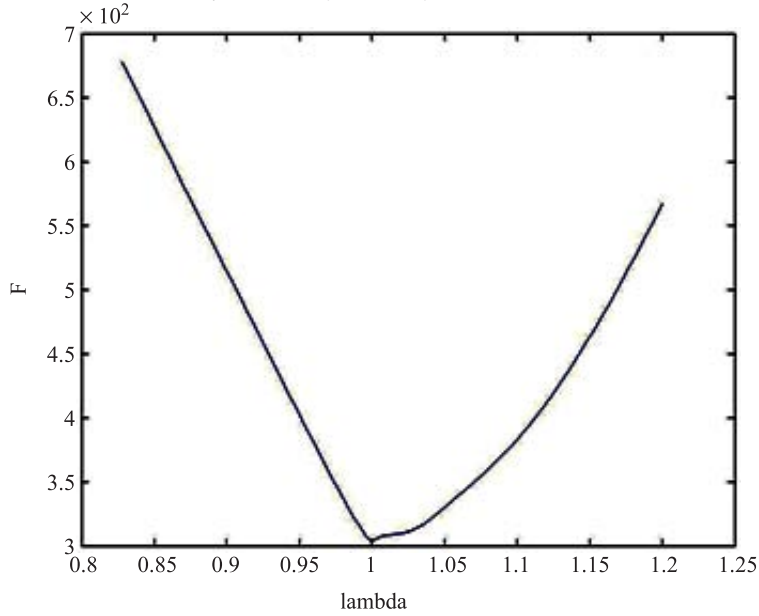


Figure 8. The cost function of the sunspot series ss_{10} , ss_{20} and ss_{40} , calculated by RLS for 200 λ values in the same range ($\lambda_c = 1$, $R = 0.4$), using the same model dimension ($na = 20$). We can see a smooth V-curve in all the cases.

Benchmark ss₉0: $\lambda_{0\text{ pt}} = 1$ $F_0\text{ pt} = 30329.1$ tiempo = 94.218



Benchmark ss₉0: $\lambda_{0\text{ pt}} = 0.999421$ $F_0\text{ pt} = 30308.9$ tiempo = 34

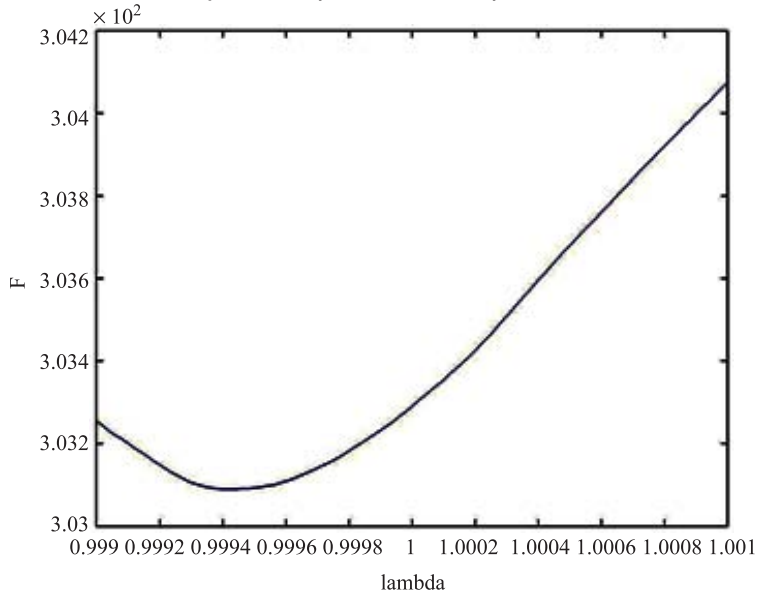


Figure 9. The cost function of ss₉₀ series calculated by RLS for 100 λ values in two different ranges ($R = 0.4$ and $R = 0.002$), both centered in $\lambda_c = 1$ and using the same model dimension ($na = 20$). We can see how the optimal λ value is not 1, and how it can be found doing the search range smaller.

Table 2. Main PARLS parameters tuned.

Parameter	Tuned value
<i>na</i>	20
λc	1
<i>R</i>	0.05
RED	2
PUN	11
PHN	4

Table 3. A summary of some results. This table shows a comparison of results between RLS search (RLS 11) and PARLS heuristics for several series with the same tuned parameters. The better algorithm is PARLS for all the series. See the conclusions section for more details.

Series	TSN	PHS	(a) RLS 11 F	(b) PARLS F	Better algorithm	PARLS vs. RLS 11
ss_00_90	36524	9131	$3.125670e + 5$	$2.792250e + 5$	PARLS	12%
ss_00_40	18262	4565	$1.561340e + 5$	$1.344990e + 5$	PARLS	16%
ss_50_90	18262	4565	$1.550950e + 5$	$1.465290e + 5$	PARLS	6%
ss_00	3652	913	$3.025500e + 4$	$1.238500e + 4$	PARLS	144%
ss_10	3652	913	$3.020200e + 4$	$1.105800e + 4$	PARLS	173%
ss_20	3653	913	$3.261300e + 4$	$1.884100e + 4$	PARLS	73%
ss_30	3652	913	$3.038700e + 4$	$1.556800e + 4$	PARLS	95%
ss_40	3653	913	$3.592300e + 4$	$2.237500e + 4$	PARLS	61%
ss_50	3652	913	$3.578300e + 4$	$2.169700e + 4$	PARLS	65%
ss_60	3653	913	$2.924200e + 4$	$2.482800e + 4$	PARLS	18%
ss_70	3652	913	$2.826900e + 4$	$2.488500e + 4$	PARLS	14%
ss_80	3653	913	$3.604300e + 4$	$3.195270e + 4$	PARLS	13%
ss_90	3652	913	$3.032900e + 4$	$2.960300e + 4$	PARLS	2%

Acknowledgement

This work has been developed thanks to TRACER project, MCYT Spain, TIC2002-04498-C05-01.

References

- Goldberg, D. E. 1989, *Genetic Algorithms in Search, Optimization and Machine Learning*. Addison Wesley.
- Gómez, J. *et al.* 2001, “Sintonización de un sistema paralelo para la Identificación de Sistemas”, In: Univ. de Extremadura (ed.): I Congreso Español de Algoritmos Evolutivos, 131–138.
- Gómez, J. *et al.* 2002, “Diseño de un coprocesador reconfigurable para un algoritmo adaptativo paralelo de Identificación de Sistemas”. In: Univ. de Granada (ed.): II Jornadas sobre Computación Reconfigurable y Aplicaciones, 221–226.
- Gómez, J., Sánchez, J., Vega, M. 2003, “Using Neural Networks in a Parallel Adaptive Algorithm for the System Identification Optimization”, *Lecture Notes in Computer Science*, 2687: Artificial Neural Nets Problem Solving Methods, 465–472.

- Kirkpatrick, S., Gelatt, C., Vecchi, M. 1983, "Optimization by Simulated Annealing", *Science*, **220**, 4598, 671–680.
- Ljung, L. 1991, System Identification Toolbox. The Math Works Inc.
- Ljung, L. 1999, System Identification – Theory for the User. Prentice-Hall, London.
- NOAA's National Geophysical Data Center (2004). Sunspot numbers.
<http://www.ngdc.noaa.gov/stp/SOLAR/ftpsunspotnumber.html>
- Rechenberg, I. 1973, Evolutionsstrategie: Optimierung technischer systeme nach prinzipien der biologischen evolution. Formmann-Holzboog Verlag.
- ROB, Royal Observatory of Belgium, Brussels (2004). Sunspot Data Series.
<http://sidc.oma.be/html/sunspot.html>
- Söderström, T. *et al.* 1989, System Identification. Prentice-Hall, London.

Preface

In the second year of my graduate studies in 1971 at the University of Texas, there was great excitement in the Department of Astronomy at the upcoming occultation of the bright star β Scorpii by Jupiter. Such occultations of very bright stars are rare and the refractive deformation of the star's brightness as it is occulted gives important information on the atmospheric structure of Jupiter. For such occultations you have to be in the right place at the right time — the path of the occultation on the ground is narrow.

A young research scientist who had only joined the staff at Texas a year before I came as a graduate student, Ed Nather, was packing up to make an expedition with his high-speed photometer to an exotic observatory on the far side of the planet to Texas, a place I first heard of then with the romantic name, *Nainital*. For reasons that remain a mystery, the name Nainital struck a chord and I knew that one day I, too, would go there. I could not know then that it would be 26 years before I did.

Over the decades since its founding in 1954 in Varanasi, then moving in 1955 to Nainital and finally to Manora Peak in 1961, the Uttar Pradesh State Observatory evolved into the State Observatory at Manora Peak, and then, in 2004, into the Aryabhata Research Institute in Observational Sciences (ARIES). ARIES has contributed significantly to the development of astronomy over its five decades and maintains a wide range of close collaboration with the rest of the Indian institutions as well as with the rest of the world. In 2004, this institute completed 50 years of scientific research, discovery and training. As part of its Golden Jubilee year celebrations we organised the **ARIES International Workshop on Asteroseismology**.

During its 50 years the diverse research interests of ARIES have been in the fields of photometric study of galaxies, stellar variability, stellar energy distribution, star clusters, planetary physics, solar activity and studies related to molecular lines in the sun. A particular research strength stemmed from that visit of Ed Nather all those years ago, since when ARIES has done high-speed photometry on many kinds of objects, with emphasis on asteroseismology — the observational study of stellar structure using surface pulsations. The clear skies and high transparency make the Nainital site excellent for this purpose, and the ARIES 1-m telescope is in high demand for asteroseismic studies.

This is particularly true for multi-site campaigns on pulsating stars from across the HR diagram: planetary nebulae nuclei, white dwarfs, subdwarf B stars, and main sequence B, A and F pulsators are some examples. For such campaigns it is critical to have continuous coverage of the observations without gaps because of the day–night cycle; for this observatories spread around the world in longitude are needed, and Nainital is sought by all, just as it was for the Jupiter occultation, because of its staff, facilities, observing conditions, and because it is in the right place. There are few observatories in good sites at the ARIES longitude. ARIES has been an

important leader and participant in many Whole Earth Telescope (WET) multi-site photometric campaigns on asteroseismic targets, and has been the leader of other such campaigns.

My own path to Nainital was circuitous. I left the US for South Africa in 1977 where I spent 24 years, and from there I made several trips to India. The first was to the International Astronomical Union (IAU) General Assembly in Delhi in 1985, when I got as close to Nainital as Corbett Park. Finally, in 1997 Ram Sagar and I, along with S. Seetha of the Indian Space Research Organisation (ISRO) and Peter Martinez of the South African Astronomical Observatory (SAAO), initiated the Nainital–Cape Survey for rapidly oscillating Ap star in the northern hemisphere. With that I began my regular trips to observe and work at ARIES.

Thus in early 2004 when Ram Sagar proposed the **ARIES International Workshop on Asteroseismology** and began assembling a Scientific Organising Committee (SOC), he asked me to act as co-chair, with the particularly pleasant duty of inviting the overseas participants. With the invaluable advice of our SOC we put together the outstanding group of scientists from India and elsewhere whose papers you will find in these proceedings. Our SOC was comprised of U. S. Chaubey (ARIES), Peter Martinez (SAAO), S. Seetha (ISRO), N. M. Ashok (Physical Research Laboratory, Ahmedabad – PRL), A. V. Raveendran (Indian Institute of Astrophysics – IIA) and K. P. Singh (Tata Institute for Fundamental Research – TIFR). We tried to limit the participation to 50 attendees, but were not completely successful, given the popularity of the meeting!

There were 48 participants from many institutions in India and 16 overseas participants from Austria, Canada, France, Japan, Korea, Portugal, Ukraine, UK, USA and Uzbekistan. Intentionally, we had a significant contingent of graduate students, giving the workshop a strong training component. The three days of the workshop were intensely packed with formal talks and informal discussions. With the meeting being held at the peaceful Uttaranchal Academy of Administration in upper Nainital, and with all participants staying there and taking all meals together, the scientific discussions and social interactions continued from morning until late into the night — in fact, for some participants until *very* late into the night!

The papers in these proceedings show the excitement and active research in asteroseismology from researchers from India and around the world. They emphasise the important role that ARIES plays in this research and they emphasise the impact that an ARIES 3-m telescope on Devastal will have, with ARIES as the leader of many international collaborations, and for the training of new generations of astronomers. During my stay at ARIES prior to this workshop I was impressed with the knowledge and unbounded enthusiasm of the graduate students there.

Socially, there were two highlights to the meeting. The most important one was the Inaugural Ceremony with the keynote speech by Prof. R. C. Pant, the Chancellor of the Kumaun University, Nainital. The Uttaranchal Song and Drama Division, Dept. of Information & Broadcasting, Govt. of India, Nainital performed regional songs and dances from the rich variety of all of India, to the great delight of the participants.

The second highlight was the Director's dinner at the Boat House Club. It was that evening standing with a group of students out on the deck of the Boat House Club looking at the magical lights and night-time stars reflected in the lake itself that I thought to myself, after so many visits to Nainital, and decades after leaving graduate school where I first heard its name, for me now it is *Texas* that is the exotic name of a faraway place on the other side of the world. That is the great pleasure of being an astronomer: we are truly a global village without borders and a global culture without barriers. The **ARIES International Workshop on Asteroseismology** expanded our scientific knowledge, our collaborations, and many new friendships were made. These proceedings record the exciting science that we discussed. The future of ARIES is bright.

Don W. Kurtz

Co-Chairman
Scientific Organizing Committee

Asteroseismology: Past, Present and Future

D. W. Kurtz

Centre for Astrophysics, University of Central Lancashire, Preston PR1 2HE, UK.

e-mail: dwkurtz@uclan.ac.uk

Abstract. Asteroseismology studies stars with a wide variety of interior and surface conditions. For two decades asteroseismic techniques have been applied to many pulsating stars across the HR diagram. Asteroseismology is now a booming field of research with stunning new discoveries; I highlight a personal selection of these in this review, many of which are discussed in more detail elsewhere in these proceedings. For many years the Nainital–Cape Survey for northern roAp stars has been running at ARIES, so I emphasise new spectroscopic results for roAp stars and point out the outstanding prospects for the planned ARIES 3-m telescope at Devastal. High precision spectroscopy has revolutionised the asteroseismic study of some types of stars – particularly solar-like oscillators and roAp stars – while photometry is still the best way to study the frequency spectra that are the basic data of asteroseismology. New telescopes, new photometers and space missions are revolutionising asteroseismic photometry. In addition to the ground-based potential of asteroseismic spectroscopy, India has the knowledge and capability for space-based asteroseismic photometry. The future for asteroseismology is bright indeed, especially for Indian astronomers.

Key words. Asteroseismology—stellar pulsation—variable stars.

1. Introduction

In 1926 in the opening paragraph of his now-classic book, *The Internal Constitution of the Stars*, Sir Arthur Eddington lamented, “What appliance can pierce through the outer layers of a star and test the conditions within?” While he considered theory to be the proper answer to that question, there is now an observational answer: asteroseismology.

In just the last few years, new observational parameter space has opened up that has revolutionised observational asteroseismology. This parameter space was pioneered in the study of extra-solar planets. The idea is simple – in fact, it is Tychonic: obtain the highest possible precision – in this case, increase the precision of radial velocity measurements by orders of magnitude. This simple idea began with observations on modest-sized telescopes in the 1-m to 4-m aperture range to bring radial velocity precision to the 1 m s^{-1} level, and better. There are now high resolution spectrographs that have been optimised for the best signal-to-noise, hence the highest precision radial velocities. Two examples are HARPS on the ESO 3.6-m telescope at La Silla, Chile (Mayor *et al.* 2003) and ESPeDOs on the Canada–France–Hawaii Telescope

(Manset & Donati 2003). The prospects for high-precision spectroscopy in the field of asteroseismology are so great that strong consideration should be given to the development and building of such a spectrograph for the planned 3-m ARIES telescope on Devastal. In the interim, some exploratory radial velocity studies on roAp stars should be started in India as an extension of the Nainital–Cape roAp star survey. I discuss this further below.

Following the successes with modest-aperture telescopes, this high-precision idea was then extended to overcome what had been long-resisted: the use of the biggest optical telescopes to look at bright stars. That idea is no revelation to solar astronomers, but especially to extra-galactic astronomers it was anathema. Use the biggest telescopes, telescopes that can see to the beginning of the universe, to look at the brightest stars in the sky? Telescope allocation committees did not take this idea seriously until recently, but with the excitement of the discovery of planets around other stars, and, hot on its heels, the discovery of solar-like oscillators, that is exactly what became not just acceptable, but highly desirable. With big telescopes, bright stars and spectrographs designed for stability, precision radial velocity measurements with standard deviations as low as of 1.9 cm s^{-1} , something approaching solar-like precision, are now being obtained. That record is for α Cen observed with the Very Large Telescope (VLT) at the European Southern Observatory (ESO) and with the Anglo-Australian Telescope (Bedding *et al.* 2004). This state-of-the-art study of solar-like oscillators now has a similar precision to that of helioseismology two decades ago. Even for magnitude 8 stars precisions of $1\text{--}2 \text{ m s}^{-1}$ are now being obtained with the largest telescopes. The new parameter space is that of simultaneous high spectral resolution, high time resolution and high signal-to-noise ratio — a triple demand that calls for plenty of photons. The regime where solar astronomers have always worked is now also the realm of stellar astronomy.

Of course, the detection of oscillations for asteroseismology and the discovery of extra-solar planets are not just using the same observing techniques. There is the possibility of confusion between the two, linking the fields intimately, and many stellar pulsators are such accurate clocks that they can be used to discover extra-solar planets. These two areas of study are thus the twin goals of new satellite missions: MOST, launched in 2003 June, and COROT and Kepler to be launched in this decade. All of these space missions are set to revolutionise asteroseismology with a factor of 10 improvement in photometric observational precision, bringing asteroseismology much closer to the current state-of-the-art in helioseismology.

Much of the excitement in asteroseismology presently is in the prospect of large frequency sets for solar-like oscillators, eventually to perform inversions and map out the interior sound speed as is done in helioseismology. I take a much more relaxed view of what constitutes asteroseismology: if a star has at least two pulsation frequencies that can be used to constrain stellar structure models, or stellar atmosphere models, then that is asteroseismology. With that definition the HR diagram is covered in asteroseismic objects to study. Asteroseismology is now a booming field of research with stunning new discoveries; I highlight a personal selection of these in this review, many of which are discussed in more detail elsewhere in these proceedings.

2. An asteroseismic HR diagram

At all current asteroseismology meetings the “pulsation HR diagram” produced by Jørgen Christensen-Dalsgaard is frequently presented to put particular classes of stars

into perspective. This was certainly true at the *ARIES International Workshop on Asteroseismology*, therefore it is reproduced here in its black-and-white version in Fig. 1 where many of the major classes of pulsating stars are noted.

2.1 *p* modes and *g* modes

There are two main sets of solutions to the equations of motion for a pulsating star, and these lead to two types of pulsation modes: *p* modes and *g* modes. For the *p* modes or pressure modes, pressure is the primary restoring force for a star perturbed from equilibrium. These *p* modes are acoustic waves and have gas motions that are primarily vertical. For the *g* modes or gravity modes, buoyancy is the restoring force and the gas motions are primarily horizontal. There are two other important properties of *p* modes and *g* modes: (1) as *n*, the radial overtone, increases, the frequencies of the *p* modes increase, but the frequencies of the *g* modes decrease; (2) the *p* modes are most sensitive to conditions in the outer part of the star, whereas *g* modes are most sensitive to the core conditions.

It is this sensitivity in particular to the core conditions that has made the discovery of *g* modes in the sun such a long-sought goal. While such a detection has been claimed for the sun repeatedly – without general acceptance yet – *g*-mode pulsators are common amongst other types of stars, even some, the γ -Dor stars, not very much hotter than the sun and overlapping with the solar-like oscillators, keeping hope alive that *g* modes may eventually be detected with confidence in the sun. There are three places in Fig. 1 where there are *p*-mode and *g*-mode pulsators of similar structure: for the β Cep (*p* mode) and Slowly Pulsating B (SPB – *g* mode) stars on the upper main sequence; for the δ Sct (*p* mode) and γ Dor (*g* mode) stars of the middle main sequence; and for the EC 14026 sub-dwarf B variables (sdBV – *p* mode; PG 1716 – *g* mode).

3. sdBV and PG 1716+426 stars

The discovery of the EC 14026 stars, the sdBV stars, was made observationally (Kilkenny *et al.* 1997) at the same time that their existence was predicted theoretically (Charpinet *et al.* 1996). The sdBV stars are low-*l*, non-radial *p*-mode pulsators with $65 \text{ s} \leq P \leq 500 \text{ s}$, and amplitudes from 1 mmag up to 0.3 mag. They are all multiperiodic with 2 to 50 modes. Driving is caused by the κ -mechanism operating on a metal line opacity bump in a layer where Fe has been radiatively levitated and concentrated.

This is one of the several classes of stars where radiative diffusion is an important astrophysical process — in this case necessary to increase the Fe concentration in the driving region to produce the observed pulsations. Such radiative diffusion is now a part of the standard solar model, but is best tested in other classes of stars: particularly the peculiar A stars, but also, as here, the sdBV stars. Hence asteroseismology provides important tests of a process now necessary to understand the sun. See Sylvain Turcotte's discussion in these proceedings, and Richard *et al.* (2004) for a recent review of diffusion.

The sdBV stars are Extreme Horizontal Branch stars that have been stripped of their H envelopes, leaving behind an approximately $0.5 M_{\odot}$ He star wrapped in a miniscule H-rich atmosphere with a mass no greater than $10^{-4} M_{\odot}$. They are all in binary systems. There is the second group of sdBV stars that pulsate with periods around one hour, the

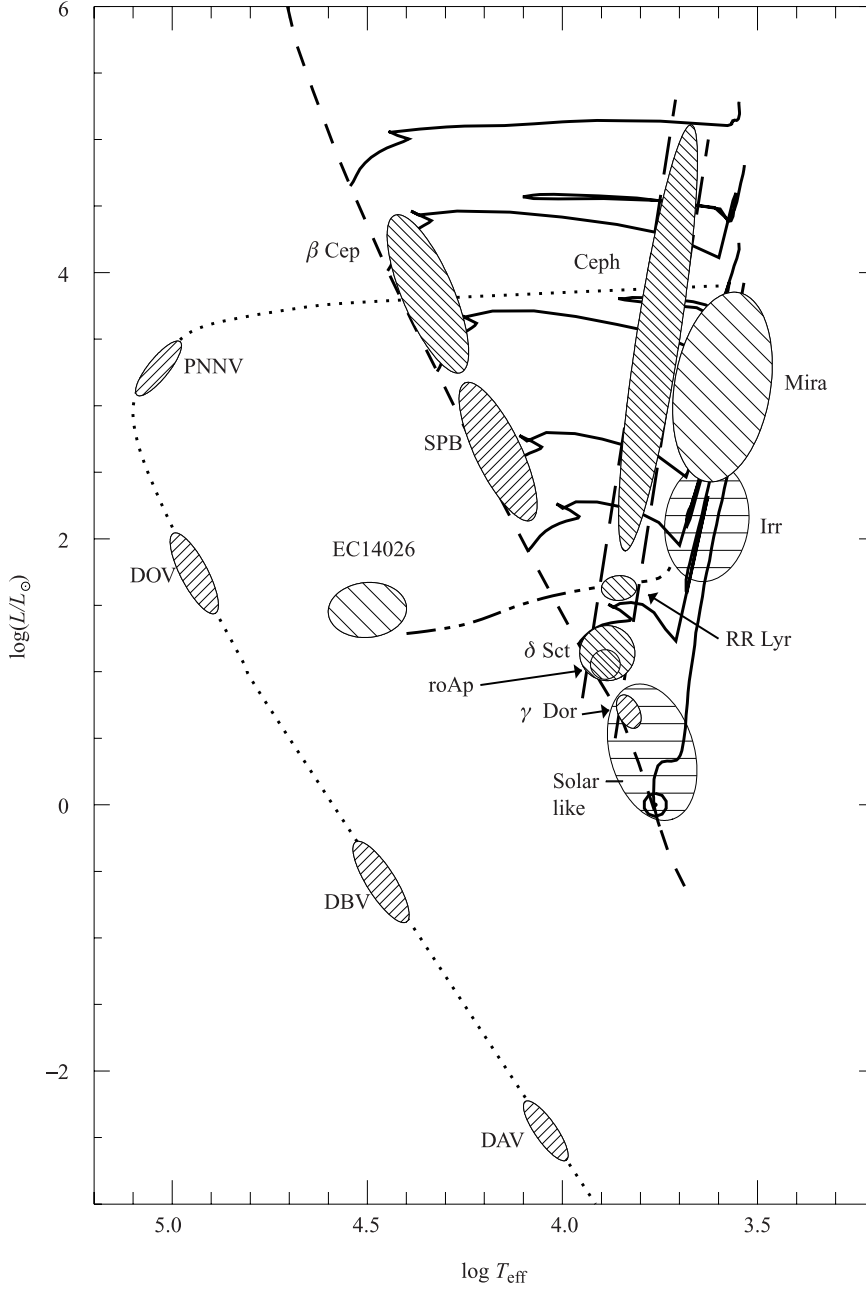


Figure 1. A pulsation HR diagram showing many classes of pulsating stars for which asteroseismology is possible. Courtesy of Jørgen Christensen-Dalsgaard.

PG 1716 + 426 stars (Green *et al.* 2003); these are g -mode pulsators that are slightly separated from the p -mode EC 14026 stars in the HR diagram. Fontaine *et al.* (2003) have found that the same κ -mechanism operating mostly on Fe opacity (in a layer with radiatively enhanced Fe abundance) drives the pulsations in these stars. Their models

indicate that the excited modes may be $l = 3$ or 4 , higher than the $l = 1$ or 2 degrees initially proposed for these stars. If correct, these higher degrees will naturally explain the low observed amplitudes in the PG 1716 + 426 stars. This field is very young, but has much promise for the study of p modes and g modes in stars of similar structure, but slightly different temperature. See Simon Jeffrey's discussion of sdBV stars in these proceedings with a demonstration of the spectacular photometric results that are coming from the UltraCam photometer. A similar, but smaller version of UltraCam would be most useful on the ARIES 1-m telescope; building such a photometer in India should be considered.

One of the most exciting of the EC 14026 stars is PG 1336–035. It is an sdBV star in an eclipsing binary with an orbital period of 2.4 hr and with an M dwarf companion of similar radius to the sdBV star. The Whole Earth Telescope (WET) devoted two campaigns to this star with extensive Indian participation (individual light curves and Fourier transforms can be viewed on the WET web site for all extended coverage campaigns). During primary eclipse as the M star passes in front of the sdBV star, the pulsation amplitude changes as parts of the non-radial modes are selectively covered. Removal of the eclipse light curve to study these amplitude changes suggests that the pulsation axis may be the tidal axis between the two stars. While there is a wealth of theoretical discussion of tidally induced and modulated stellar pulsation, there have been few observations to test the theory, giving PG 1336 great promise for that.

Another exciting sdBV star, the subject of WET Xcov23 (the 23rd WET extended coverage campaign) is KPD 1930 + 2752, an ellipsoidal variable with over 44 periods, and an orbital period of 2.28 hr. The companion is unseen and exceeds the mass of KPD 1930 + 2752. In fact, the projected orbital velocity of 349 km s^{-1} implies a total mass for the system of *at least* $1.47 M_{\odot}$ — greater than the Chandrasekhar limit! Gravitational radiation is expected to cause the system to merge in about 200 Myr making KPD 1930 + 2752 the first good candidate of a Type Ia supernova (SNe Ia) double-degenerate progenitor.

For application of asteroseismology, it is necessary to have clear mode identifications of n , l and m for each mode. The combination of line profile and radial velocity studies from spectroscopy and concurrent photometry is best for this. Now in addition to the photometric WET there is the Multi-site Spectroscopic Survey Telescope (MSST). A large campaign by the MSST in 2002 was awarded 113 nights of spectroscopic time to observe the sdBV star PG 1605 + 072 in conjunction with 127 hr of photometric observations obtained by WET in Xcov22. First results for the spectroscopy can be found in O'Toole *et al.* (2004). The technique of using multisite campaigns for both spectroscopy and photometry is very fruitful and will be the standard for the best observational studies to come.

4. β Cep and SPB stars

The β Cep stars are (primarily) p -mode main sequence pulsators of spectral type B0–B2. They have large convective cores which make them interesting targets for asteroseismology, because of the uncertainty of modelling such cores. Although they are multi-periodic, they are not known to have large numbers of pulsation modes such as are found for solar-like or white dwarf oscillators. Nevertheless, much can be learned from just a few modes, and there have been major recent asteroseismic successes for some β Cep stars.

Aerts *et al.* (2003) studied 21 yr of multi-colour photometry for the β Cep star HD 129929 (V836 Cen). They identified an $l = 1$ triplet, a doublet they attributed to two m -modes of an $l = 2$ quintuplet and an $l = 0$ radial mode, all with frequencies in the range of $6\text{--}7\text{ d}^{-1}$. The models that best fitted the observed frequencies had a core overshooting parameter of $\alpha_{\text{over}} = 0.1$ and non-rigid rotation, with the core rotating four times faster than the surface. If this result stands the test of time, it is the first detection of differential rotation with depth in a star other than the sun.

For mode identification De Ridder *et al.* (2004) and Aerts *et al.* (2004) have obtained extensive photometric and spectroscopic data sets for the β Cep star ν Eri: 600 hr of differential photometry on 148 nights and 2442 spectra for line profile and radial velocity studies obtained on 118 nights. The frequencies are in the $5\text{--}8\text{ d}^{-1}$ range and there is some suggestion from the lowest frequencies of the presence of g modes, as well as p modes. If this is confirmed, it will make ν Eri an SPB star, as well as a β Cep star.

As can be seen in Fig. 1, the g -mode SPB stars are somewhat cooler than the p -mode β Cep stars. For decades the driving mechanism for the β Cep stars was a mystery. Then the new opacities increased the driving from the metal lines and Fe was found to be the primary source of the opacity for the κ -mechanism for both the β Cep stars and for the SPB stars. The agreement between the theoretical instability strips for these two classes and the observations is striking, particularly in the case of the SPB stars where the observations that fill the instability strip came after the theoretical calculation had been made (see Fig. 1 of Waelkens *et al.* 1998).

Given that the κ -mechanism is operating on Fe and the Fe-peak elements to drive the pulsations in β Cep stars, it was supposed that the lower metallicity of the Magellanic Clouds would mean that these stars would be rare there. Kołaczkowski *et al.* (2004) have destroyed that idea with the amazing discovery of 64 β Cep stars in the LMC. Their periods are much longer than those found for β Cep stars in the galaxy, suggesting that they are pulsating in low-overtone g modes. This means that β Cep stars can pulsate in both p modes and g modes, depending on their metallicity. The existence of the LMC β Cep stars also suggests that radiative diffusion is operating to enhance the Fe in the driving region, much as in the sdBV stars, so here is another case of asteroseismic targets for which radiative diffusion is an important process. Kołaczkowski *et al.* also found multiperiodic SPB stars in the LMC giving great promise for further studies of these two related classes of stars in a metallicity environment very different from our own galaxy.

5. δ Sct and γ Dor stars

Before the recent discovery of the solar-type oscillators the δ Sct stars (p -mode pulsators with periods between about 20 min and 6 hr) were the only non-degenerate stars with many pulsation frequencies offering a lot of scope for asteroseismology. The problem currently for asteroseismology of these stars is that there are hundreds of excited modes predicted theoretically, while at most only a few dozen are observed, so that modes cannot be identified reliably and constraints on models are weak. It is rotation, and particularly differential rotation, that generates a vast number of possible pulsation frequencies in the models. There is also the problem of mixed modes – modes that behave as g modes in the interior but as p modes in the envelope.

Again, as can be seen in Fig. 1, there is the region of g -mode pulsation for the γ Dor stars just cooler than the primarily p -mode pulsation of the δ Sct stars. Unlike the β Cep and SPB stars, though, in this case the driving is probably not by the same mechanism. The δ Sct stars are driven by the κ -mechanism operating mostly on He II. The best models for driving in the γ Dor stars suggest that driving mechanism is convective blocking of radiation in stars where the convective zone reaches down to 200 000–400 000 K. Recently Warner *et al.* (2003) have produced a theoretical instability strip for the γ Dor stars that matches well 30 known stars of the class. The δ Sct stars and γ Dor stars are discussed in detail in these proceedings by Marie-Jo Goupil and Gerald Handler, respectively. Reza Samadi makes predictions in these proceedings that suggest it may be possible to find δ Sct stars and γ Dor stars that are also solar-like oscillators. A project to search for such objects has been initiated by Handler as a result of this ARIES workshop.

6. Rapidly oscillating Ap (roAp) stars

The rapidly oscillating Ap (roAp) stars are chemically peculiar, strongly magnetic A stars that pulsate in high overtone acoustic modes with periods between 5 and 21 min. They have been observed intensively photometrically since their discovery (Kurtz 1982), culminating in a three-week Whole Earth Telescope campaign on the roAp star HR 1217 (Xcov20) that reached a photometric precision of $14 \mu\text{mag}$ (Kurtz *et al.* 2005). Frequency analyses of their light curves have yielded rich asteroseismic information on the degrees of the pulsation modes, distortion of the modes from normal modes, magnetic geometries, and luminosities. The latter, in particular, are derived asteroseismically and agree well with Hipparcos luminosities (Matthews *et al.* 1999).

ARIES participated in the Xcov20 of HR 1217, one of the best-studied rapidly oscillating Ap (roAp) stars, with a frequency spectrum of alternating even- and odd- l modes that are distorted by the presence of a strong, global magnetic field. Several recent theoretical studies have found that within the observable atmospheres of roAp stars the pulsation modes are magneto-acoustic with significant frequency perturbations that are cyclic with increasing frequency. To test these theories Xcov20 obtained 342 hr of Johnson B data at 10 s time resolution for HR 1217 over 35 d with a 34% duty cycle in 2000 November – December. The precision of the derived amplitudes is $14 \mu\text{mag}$, making this one of the highest precision ground-based photometric studies ever undertaken. Substantial support has been found for the new theories of the interaction of pulsation with the strong magnetic field. In particular, the frequency jump expected as the magnetic and acoustic components cycle through 2π radians in phase has been found. Additionally, comparison of the new 2000 data with an earlier 1986 multi-site study shows clear amplitude modulation for some modes between 1986 and 2000. The unique geometry of the roAp stars allows their pulsation modes to be viewed from varying aspects with rotation, yielding mode identification information in the rotational sidelobes that is available for no other type of pulsating star. Those rotational sidelobes in HR 1217 confirm that two of the modes are dipolar, or close to dipolar; based on the frequency spacings and Hipparcos parallax three other modes must be either $l = 0$ or $l = 2$ modes, either distorted by the magnetic field, or a mix of m -modes of given l where the mixture is the result of magnetic and rotational effects. Elsewhere in these proceedings, S. Seetha discusses the Indian contribution to this WET campaign and others.

The Nainital–Cape Survey for Northern roAp stars has been running at ARIES for some years now (see the discussions by Ram Sagar and by S. Joshi in these proceedings), and has discovered one roAp star, HD 12098 which is of great interest, as it has the second-highest surface magnetic field strength, ~ 6.5 kG, of any roAp star (Ryabchikova *et al.* 2004). It is also the only far northern declination roAp star known. Outstanding results for this star are presented in these proceedings by V. Girish from a multi-site campaign led by him and S. Seetha, in a follow-up to the Nainital–Cape Survey discovery.

Although the survey has found only one new roAp star so far, many other northern searches for roAp stars have found *none*. The low photometric amplitude of these stars makes finding them and studying them photometrically a real challenge. It has now been found that much higher signal-to-noise can be obtained spectroscopically – at least with 8-m class telescopes – while the oscillations may be discovered more efficiently spectroscopically with telescopes in the 3-m class with optimised spectrographs, than they can be photometrically. I illustrate some exciting new results for a few of these stars for which we have VLT spectra at high time resolution.

Because of the high overtones, the vertical wavelengths of the pulsation modes in roAp stars are short compared to atmospheric continuum optical depth $\tau_{5000} = 1$, so that the pulsation behaviour can be probed as a function of atmospheric depth by examining lines of different strengths that form at different depths, and by studying pulsation amplitude and phase within the line-forming layer of a single ion by taking cuts through a single spectral line. Recent analyses also show that the atmospheres of these stars are chemically highly stratified, accentuating the depth sensitivity obtained by studying individual lines of different ions.

The spectra of many roAp stars show a strong core-wing anomaly in the hydrogen lines, particularly the H α line (Cowley *et al.* 2001; Kochukhov *et al.* 2002). This indicates abnormal atmospheric structure, as does the fact that consistent abundances for the first and second ionisation states of Rare Earth elements (REE), particularly Nd and Pr, cannot be found for these stars (Ryabchikova *et al.* 2002) with standard A star atmospheric models.

In general, for the roAp stars, Fe lines originate near a radial node around optical depth $\tau_{5000} = 1$ with little, or no radial velocity variations seen; the core of the H α line forms higher in the atmosphere and shows radial velocity variations with amplitudes up to 5 km s^{-1} (10 km s^{-1} peak-to-peak); lines from the first and second ionisation states of the REE Pr and Nd arise in layer above optical depth $\tau_{5000} = 10^{-3}$ and show amplitudes as large as 5 km s^{-1} .

The high-amplitude record holder is HD 99563, discovered to be an roAp star by Dorokhova & Dorokhov (1998; see also Tanya Dorokhova’s discussion in these proceedings). Handler *et al.* (2002) showed a range from essentially zero amplitude up to 6 mmag, most probably caused by rotation of star which they find to be the shortest of any roAp star, $P_{\text{rot}} = 1.456$ d. Elkin *et al.* (2005) have studied a VLT data set for this star and find remarkable results.

For example, while the amplitude measured for the whole of the H α core is 2.6 km s^{-1} , this is not the most informative way to examine the H α line variations. It is much more informative to measure the radial velocities of the line bisector at different depths in the line; this gives the distribution of the velocity field as a function of line depth, hence as a function of height in the atmosphere. This technique of studying the radial velocities as a function of depth for the line bisector is powerful for resolving

the pulsations and atmospheric physics as a function of atmospheric height; it also shows that it is not possible to study roAp stars by characterising their line profiles with any simple spherical harmonic sum of modes in a single line-forming layer, as is always done for other types of pulsating stars. Figure 2 shows the amplitude spectrum for a cut through the $H\alpha$ line at an intensity depth of 0.32 where the amplitude reaches a maximum of nearly 5 km s^{-1} . As can be seen, the S/N ratio is outstanding.

Figure 3 shows the change of amplitude and phase as a function of line depth for the narrow $H\alpha$ core. The bisector measurements in the core were made from intensity 0.27

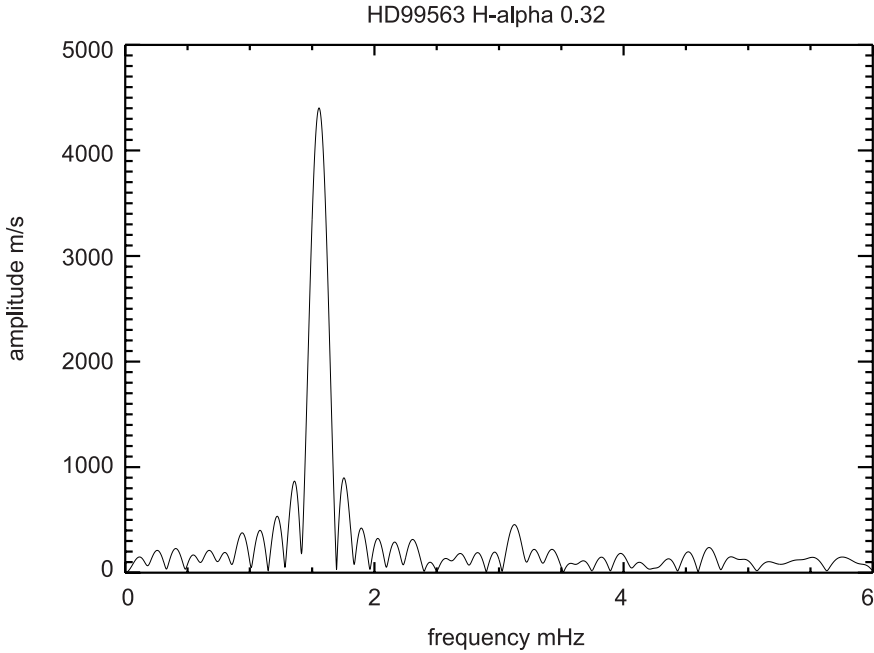


Figure 2. An amplitude spectrum for a cut through the $H\alpha$ line of HD 99563 at an intensity depth of 0.32, showing the highest amplitude in the $H\alpha$ line-forming layer. The harmonic is visible, also.

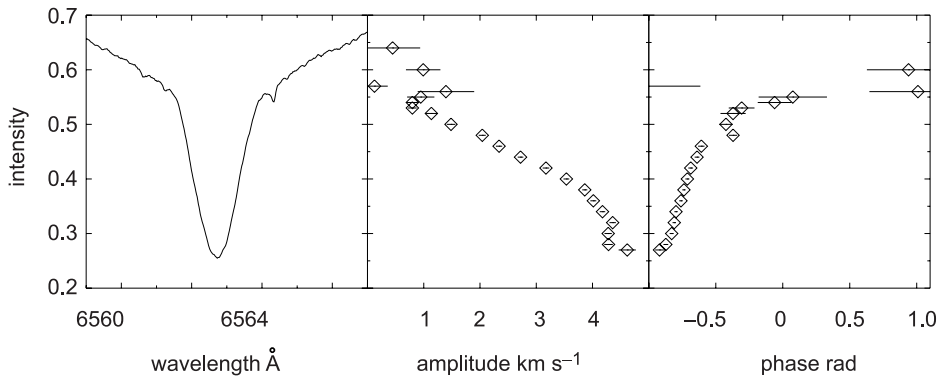


Figure 3. Pulsation amplitude and phase as a function of line depth (atmospheric height) in the $H\alpha$ core showing the dramatic changes.

to 0.54; they have good precision and show a very clear picture: the amplitude increases dramatically from 0.8 km s^{-1} to 4.5 km s^{-1} from the top to the bottom of the core, *i.e.*, with increasing atmospheric height. The formation height of the top of the narrow line core is near a radial pulsation node, and there is a hint that the bottom of the core forms near an anti-node. The phase also changes smoothly for this part of the profile. A smaller increase in the amplitude of radial velocity variations with depth in the $\text{H}\alpha$ core has also been detected for γ Equ (Sachkov *et al.* 2004), whereas no clear variation with line depth in $\text{H}\alpha$ was found for three pulsation modes in HD 166473 (Kurtz *et al.* 2003).

These atmospheric height effects are further demonstrated with VLT observations of 33 Lib (HD 137949; Kurtz *et al.* 2005). The largest amplitudes found for the principal pulsation frequency in 33 Lib are for $\text{H}\alpha$ and some lines of singly-ionised rare earth elements, Nd II, Ce II and Sm II with semi-amplitudes of $200\text{--}300 \text{ m s}^{-1}$. Lines of Nd III, Pr III and Eu II typically have amplitudes in the $120\text{--}200 \text{ m s}^{-1}$ range, while lines of Fe, Cr and Ca usually have amplitudes less than 40 m s^{-1} . Figure 4 shows amplitude spectra for the Nd II line at $\lambda 5319.815 \text{ \AA}$ and the Nd III line at $\lambda 6145.070 \text{ \AA}$ showing three of these frequencies, and for the Fe I line at $\lambda 5434.524 \text{ \AA}$ where only the principal frequency can be seen. The amplitudes of the 2.015-mHz and 4.030-mHz peaks show their independent behaviour as a function of atmospheric depth, since these three lines arise in different layers.

This, then, begins to show the power of spectroscopy compared to photometry for surveying for new roAp stars. Photometric observations have traditionally been done in the Johnson *B* filter for roAp stars, although there are observational studies in *UBVRI* and *uvby* for some stars. For 33 Lib Kurtz (1991) observed in Johnson *U* and found the principal pulsation frequency at $\nu_1 = 2.015 \text{ mHz}$ and its harmonic, with a hint that there might be a secondary frequency at 1.975 mHz , but no indication of any variation at the $\nu_2 = 1.769 \text{ mHz}$ found in the radial velocity variations and seen in Fig. 4. Figure 5 shows the amplitude spectrum for an 8.4-hr light curve obtained by Kurtz (1991) illustrating the absence of ν_2 , and the small amplitude of the harmonic.

Of course, the photometric observations (obtained in 1987) and the spectroscopic observations (obtained in 2004) are separated by 17 yr, and roAp stars are known to have mode lifetimes shorter than this in some cases. For example, Kreidl *et al.* (1991) found a new pulsation frequency for the roAp star HD 217522 in data obtained in 1989 that was definitely not present at high precision in data obtained in 1982. Because of this same possibility for 33 Lib, we obtained a new light curve for this star 75 d after the spectroscopic data were obtained. While this is not simultaneous, it is close in time for this star which has shown no large changes to its frequency spectrum over two decades of photometric observations.

Therefore the photometry is constrained to observe at one atmospheric level for one particular filter, and there may, or may not, be substantial pulsation amplitude at that level. In many cases the photometry may be probing a layer where there is a pulsation node. Figure 6 shows such a node for 33 Lib with the amplitudes and phases determined by least squares for the 2.015-mHz mode for Nd II and Nd III. For each of these ions the phases do not appear to vary as a function of atmospheric depth; those phases differ between the ions by 2.65 ± 0.31 radians, nearly in antiphase (π radians) as found by Mkrtichian *et al.* (2003). There is also a clear increase in amplitude with atmospheric height for Nd III and increase in amplitude with atmospheric depth for Nd II. This is completely consistent with Mkrtichian *et al.*'s suggestion that the lines

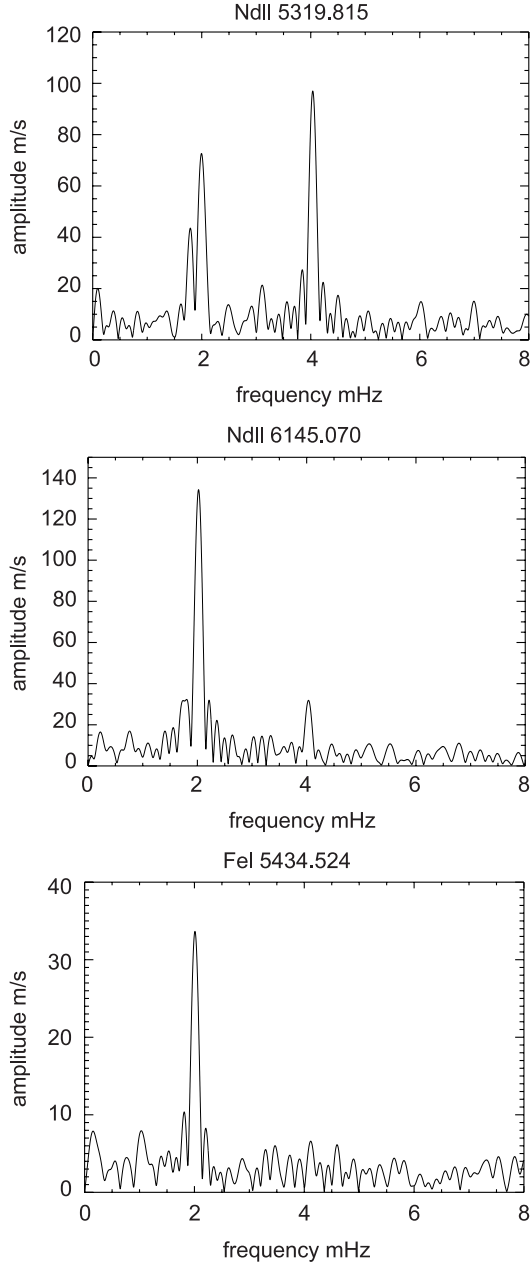


Figure 4. Top panel: the amplitude spectrum of the radial velocity variations for the Nd II $\lambda 5319.815$ Å line, the Nd III $\lambda 6145.070$ Å line and the Fe I $\lambda 5434.524$ Å line, showing variations of pulsation amplitude with atmospheric height in 33 Lib. Note the difference of ordinate scales in the panels.

of these two ions form on opposite sides of a pulsation node with Nd III above, and Nd II below the node. (See, also, Mkrtichian's discussion in these proceedings of the most peculiar star of all, the roAp star HD 101065.)

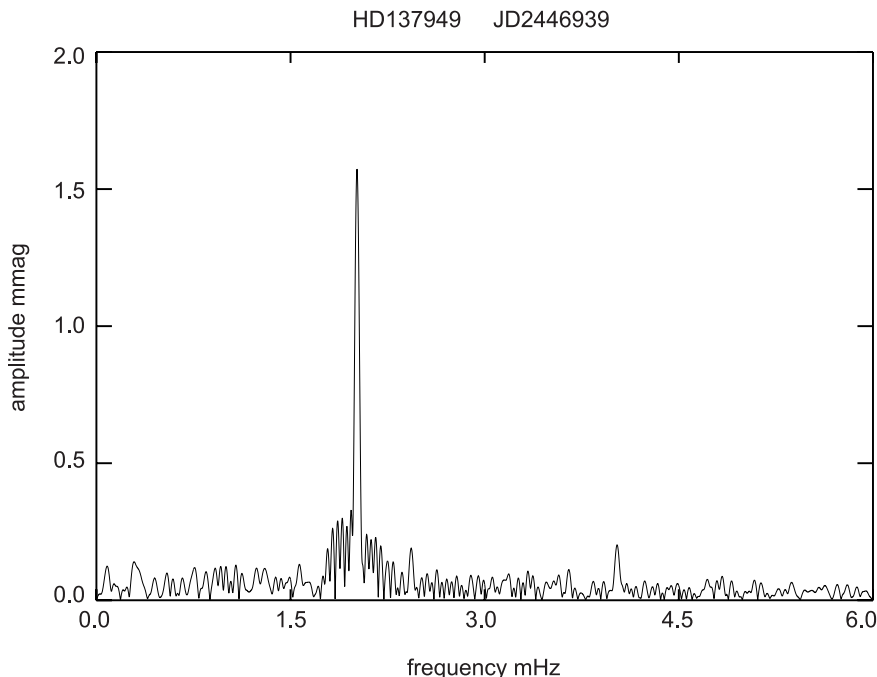


Figure 5. The amplitude spectrum of the *U* photometric variations of 33 Lib for an 8.4 hr light curve (Kurtz 1991) showing the principal frequency and its (small) harmonic. There is no evidence of $\nu_2 = 1.769$ mHz.

While the photometry may probe only one layer, the spectroscopy can pick and choose among layers by looking at lines formed at different depths and by taking cuts through lines. It is thus potentially much more sensitive to detecting and studying new variables, as long as sufficient S/N can be obtained. From the large radial velocity pulsation amplitude of HD 99563 shown in Figs. 2 and 3, it should be possible to discover some new roAp stars with telescopes in the 2-m class; tests in India should be made to examine the feasibility of this, with longer-term plans for a fuller survey for northern roAp stars with the 3-m ARIES telescope on Devastal.

One final demonstration of the sensitivity for roAp stars of the radial velocity studies is the discovery of 21-min pulsations in the luminous Ap star HD 116114 (Elkin *et al.* 2005) using high resolution spectra obtained with the UVES spectrograph at the ESO VLT. The highest amplitudes of the radial velocity variations are between 50 and 100 m s^{-1} visible in the Eu II lines. The spectral lines of La II and the core of the H α line have amplitudes of about 30 m s^{-1} . According to Cunha (2002; also see her discussion in these proceedings) it is expected that roAp stars within the instability strip will have oscillation frequencies between 2.1 and 3.2 mHz near the Zero Age Main Sequence and from 0.7 to 1.0 mHz as they near core hydrogen exhaustion. Cunha (2002) notes that the oscillations in more evolved and/or more luminous stars, having lower frequencies, could be missed more often than other, higher frequency roAp stars due to higher noise levels. The frequency obtained for HD 116114 is in good agreement with these theoretical predictions of longer-period, evolved roAp stars. The distinction in luminosity between the roAp and noAp stars, and the suggestion that

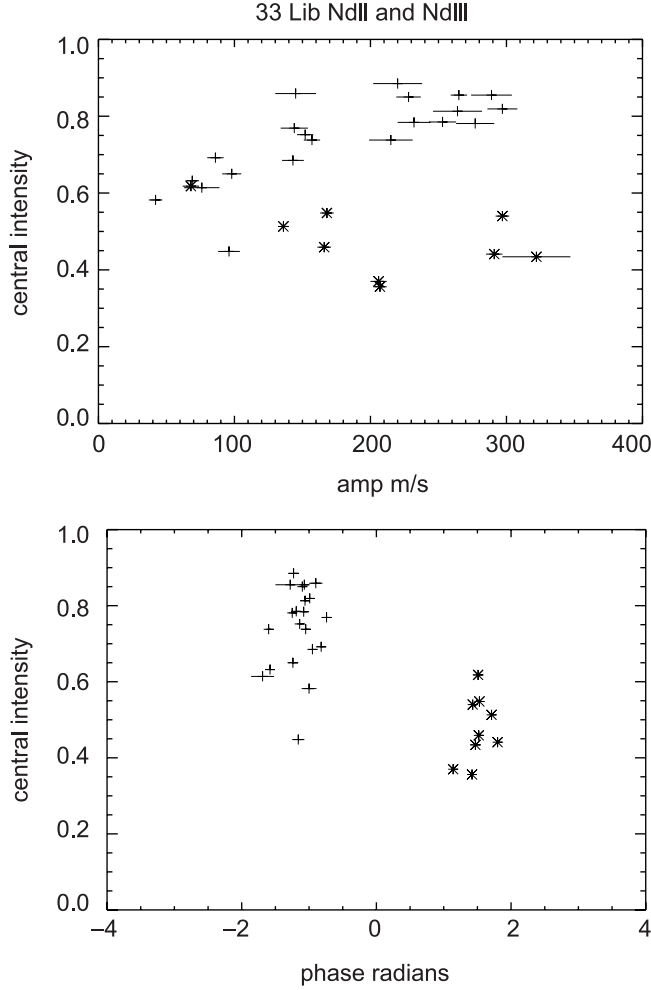


Figure 6. The amplitudes (top panel) and phases (bottom panel) of the Nd II and Nd III lines as a function of central line intensity normalised to 1 at the continuum. The smaller the value of the central intensity, the deeper the line, the higher the opacity, hence the higher in the atmospheric layer for that ion the line forms. The top panel shows the amplitudes for the Nd II lines (plus signs) and Nd III lines (asterisks). The error bars are 1σ and in some cases are smaller than the points. Because of decreasing density and ionisation equilibrium, the Nd III lines form higher in the atmosphere than the Nd II lines. The two layers lie on opposite sides of a radial pulsation node with the Nd III amplitudes increasing outwards away from the node, and the Nd II amplitudes increasing inwards away from the node.

in all roAp stars the abundance of the second ions of Pr and Nd, relatively to the abundance of the first ions, is anomalously high, needs to be revised in the light of this discovery.

Figure 7 shows some amplitude spectra for HD 116114 demonstrating the presence of the 21-min pulsations at low amplitude. Photometric searches for pulsation in this star have been unsuccessful – not surprisingly given the low radial velocity amplitude. It seems possible that many of the stars in the Nainital–Cape Survey for which we have found no photometric variations may, in the future, be discovered to be radial

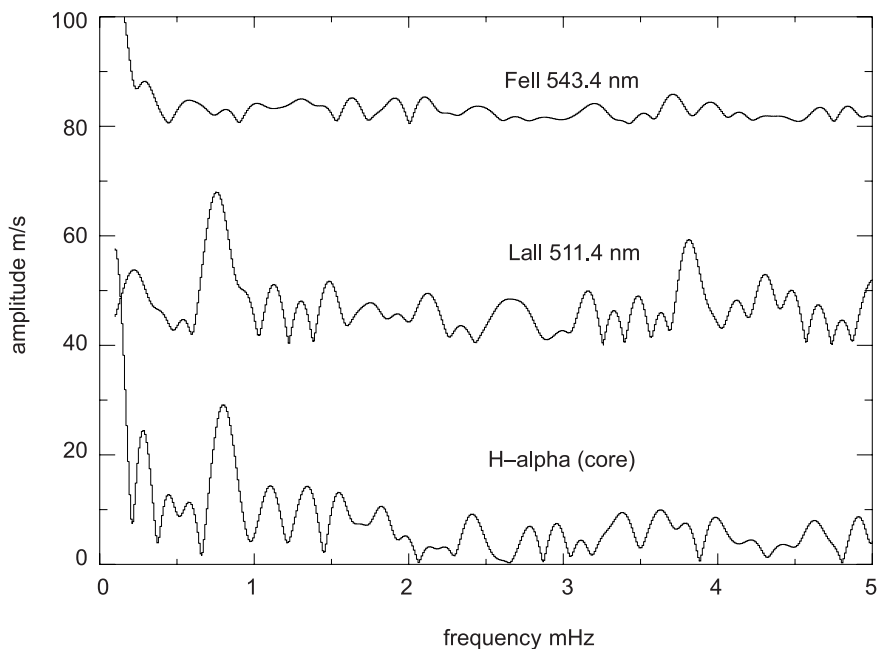


Figure 7. Amplitude spectra for HD 116114 for the sum of 11 lines and for the central component of the Eu II line at $6437 \text{ \AA}$ which show peaks near 0.79 mHz ($P = 21 \text{ min}$), and for line Cr II $5334 \text{ \AA}$ which does not show rapid oscillations. The spectra for Eu II and Cr II are shifted at 40 and 80 m s^{-1} respectively on y-axis.

velocity variables. This gives further strong support to the ARIES plans for the 3-m Devastal Telescope.

7. White dwarf variables

White dwarf variables are g -mode pulsators and are the current champions of asteroseismology. They have more frequencies detected than any other type of pulsating star, other than the sun, and theory has been more successful in extracting astrophysical information for them than for any other kind of pulsator. As can be seen in Fig. 1 there are three main regions of white dwarf pulsation: The DOV, DBV and DAV stars, where the nomenclature is D = white dwarf; V = pulsating variable; and O, B and A refer to the spectra that resemble O, B and A stars in the presence of He and H lines.

The best studies of pulsating white dwarfs have been carried out by the WET. The WET website contains a wealth of information and references to published papers from many extended coverage (Xcov) campaigns. An outstanding example is their study of the DOV star PG 1159-035 (Winget *et al.* 1991) where they found 101 independent pulsations modes. Models yielded a mass of $M = 0.586 \pm 0.003 M_{\odot}$; independent tests using distances determined from parallaxes and the mass-radius relation indicate that the quoted precision is probably correct. The periods in PG 1159 are in the range $385 \leq P \leq 1000 \text{ s}$; they are high-overtone ($n \gg l$) g -mode pulsations, as is the case for other pulsating white dwarfs. Asymptotic theory gives a clear prediction of period

spacing for such stars, and deviations from that are used to derive the compositional stratification in their atmospheres, *i.e.*, the mass of the surface He and/or H layers – possibly even resolving He³ and He⁴ layers (Wolff *et al.* 2002). PG 1159 clearly shows $l = 1$ and 2 modes, but not $l = 3$. The magnetic field strength is less than 6000 G — a very small value for a white dwarf star where field strengths are often MG. Clearly asteroseismology of white dwarf stars is highly successful in extracting astrophysically interesting information.

One of the more striking properties of white dwarfs being studied by asteroseismology now is the C/O interior composition and the potential crystallisation of the core in the most massive of the DAV stars, such as BPM 37093 (Montgomery *et al.* 1999), into earth-sized “diamonds”. These new kinds of “diamonds” are C and O in a partial to complete crystalline state of degenerate matter; this state of matter has been (briefly) produced experimentally using petawatt lasers at Lawrence Livermore Laboratory.

The structure and composition of the C/O cores of DBV models have been probed by Metcalfe *et al.* (2003). Metcalfe (2003; see, also, his discussion elsewhere in these proceedings) has even used seismic models of two white dwarfs to measure the $^{12}\text{C}(\alpha, \gamma)^{16}\text{O}$ reaction rates, finding a value consistent with laboratory reaction rates.

Perhaps even more interestingly, white dwarfs produce much of their total radiation in neutrinos; for the very hot white dwarfs the neutrino flux exceeds the photon flux. Some of the neutrinos are brehmsstrahlung neutrino pairs from electrons, but most are plasmon neutrinos that are created by photons that have an effective rest mass when travelling through the dense plasma in the white dwarf and can decay into neutrino-antineutrino pairs. Thus the cooling rate for white dwarfs is a measure of the neutrino generation rate and asteroseismology can test the cooling rate. This field is young; see O’Brien & Kawaler (2000) for a first attempt at a test of standard lepton theory in a dense plasma from asteroseismology of the DOV star PG 0122+200.

8. Conclusions

Asteroseismology has come a long way in the last two decades. At its highest current precision radial velocity studies of solar-like stars, especially α Cen are now as precise as solar observations were 20 years ago. The papers in these proceedings show the strong Indian involvement in many asteroseismic programmes on many types of stars. In particular, India is a regional leader in this field with excellent prospects for multi-site collaboration with other Asian countries, especially in photometric campaigns. Observationally, the longitude of India and its regional collaborators make this region critical to all global observational helioseismic and asteroseismic studies. See the discussions in these proceedings by Tatyana Dhorokova and Alisher Hojaev about neighbouring countries’ potential for participation with India in multi-site campaigns.

The remarkable advancement in precision of radial velocity studies with 8-m class telescopes, and now optimised spectrographs on 3–4-m class telescopes strongly supports the ARIES plan for a 3-m telescope on Devastal. This telescope will be at the forefront of many asteroseismic studies.

India’s space capabilities are impressive. With the strong collaboration between ISRO and ARIES in asteroseismology, and with the revolutionary prospects of asteroseismic satellites, it is realistic to imagine that India may build and launch such a satellite in the not-too-distant future.

The prospects look bright for asteroseismology in India. We can look forward to discussions of the developments at future ARIES International Workshops on Asteroseismology.

References

- Aerts, C. *et al.* 2003, *Science*, **300**, 1926.
Aerts, C. *et al.* 2004, *MNRAS*, **347**, 463.
Bedding, T. R., Kjeldsen, H., Butler, R. P., McCarthy, C., Marcy, G. W., O'Toole, S. J., Tinney, C. G., Wright, J. T. 2004, *APJ*, **614**, 380.
Charpinet, S., Fontaine, G., Brassard, P., Dorman, B. 1996, *ApJL*, **471**, L103.
Cowley, C. R., Hubrig, S., Ryabchikova, T. A., Mathys, G., Piskunov, N., Mittermayer, P. 2001, *A&A*, **367**, 939.
Cunha, M. S. 2002, *MNRAS*, **333**, 47.
De Ridder, J. *et al.* 2004, *MNRAS*, **351**, 324.
Dorokhova, T. N., Dorokhov, N. I. 1998, *Contributions of the Astronomical Observatory Skalnaté Pleso*, **27**, 338.
Elkin, V. G., Kurtz, D. W., Mathys, G. 2005, *MNRAS*, submitted.
Elkin, V. G., Riley, J. D., Cunha, M. S., Kurtz, D. W., Mathys, G. 2005, *MNRAS*, **358**, 665.
Fontaine, G., Brassard, P., Charpinet, S., Green, E. M., Chayer, P., Billères, M., Randall, S. K. 2003, *ApJ*, **597**, 518.
Green, E. M. *et al.* 2003, *ApJL*, **583**, L31.
Handler, G. *et al.* 2002, In: *ASP Conf. Ser.*, 256: *Observational Aspects of Pulsating B- and A Stars*, **256**, 109.
Kilkenny, D., Koen, C., O'Donoghue, D., Stobie, R. S. 1997, *MNRAS*, **285**, 640.
Kochukhov, O., Bagnulo, S., Barklem, P. S. 2002, *ApJL*, **578**, L75.
Kołaczkowski, Z., Pigulski, A., Soszyński, I., Udalski, A., Szymański, M., Kubiak, M., Zeburń, K., Pietrzyński, G., Woźniak, P. R., Szewczyk, O., Wyrzykowski, Ł. 2004, In: *ASP Conf. Ser.*, *Variable star in the Local Group* (ASP: San Francisco), **310**, 225.
Kreidl, T. J., Kurtz, D. W., Bus, S. J., Kuschnig, R., Birch, P. B., Candy, M. P., Weiss, W. W. 1991, *MNRAS*, **250**, 477.
Kurtz, D. W. 1982, *MNRAS*, **200**, 807.
Kurtz, D. W. 1991, *MNRAS*, **249**, 468.
Kurtz, D. W. *et al.* 2005, *MNRAS*, **358**, 651.
Kurtz, D. W., Elkin, V. G., Mathys, G. 2003, *MNRAS*, **343**, L5.
Kurtz, D. W., Elkin, V. G., Mathys, G. 2005, *MNRAS*, **358**, L6.
Manset, N., Donati, J. 2003, *Proceedings SPIE*, **4843**, 425.
Metcalf, T. S. 2003, *ApJL*, **587**, L43.
Metcalf, T. S., Montgomery, M. H., Kawaler, S. D. 2003, *MNRAS*, **344**, L88.
Mkrichian, D. E., Hatzes, A. P., Kanaan, A. 2003, *MNRAS*, **345**, 781.
Matthews, J. M., Kurtz, D. W., Martinez, P. 1999, *ApJ*, **511**, 422.
Mayor, M. *et al.* 2003, *The Messenger*, **114**, 20.
Montgomery, M. H., Klumpe, E. W., Winget, D. E., Wood, M. A. 1999, *ApJ*, **525**, 482.
O'Brien, M. S., Kawaler, S. D. 2000, *ApJ*, **539**, 372.
O'Toole, S. J., Falter, S., Heber, U., Jeffery, C. S., Dreizler, S., Schuh, S. L. & The MSST + Wet Teams, 2004, *Ap&SS*, **291**, 457.
Richard, O., Théado, S., Vauclair, S. 2004, *Solar Phys.*, **220**, 243.
Ryabchikova, T. *et al.* 2002, *A&A*, **384**, 545.
Ryabchikova, T., Nesvácil, N., Weiss, W. W., Kochukhov, O., Stütz, C. 2004, *A&A*, **423**, 705.
Sachkov, M., Ryabchikova, T., Kochukhov, O., Weiss, W. W., Reegen, P., Landstreet, J. D. 2004, In: *ASP Conf. Ser.* 310: *IAU Colloq. 193: Variable Stars in the Local Group*, **310**, 208.
Waelkens, C., Aerts, C., Kestens, E., Grenon, M., Eyer, L. 1998, *A&A*, **330**, 215.
Warner, P. B., Kaye, A. B., Guzik, J. A. 2003, *ApJ*, **593**, 1049.
Winget, D. E. *et al.* 1991, *ApJ*, **378**, 326.
Wolff, B., Koester, D., Montgomery, M. H., Winget, D. E. 2002, *A&A*, **388**, 320.

Theory of Asteroseismology

Hiromoto Shibahashi

Department of Astronomy, University of Tokyo, Tokyo 113 0033, Japan.

e-mail: shibahashi@astron.s.u-tokyo.ac.jp

Abstract. The increasing number of pulsating modes in a star leads to the increase of the amount of information derived from the pulsation. Probing the invisible interior of stars is a dream come true. In this review, I summarize the fundamental properties of stellar pulsations and the excitation mechanisms, emphasizing several cases of recent progress.

Key words. Asteroseismology—stars: pulsation—pulsation: excitation.

1. Introduction

We have recently witnessed a new development in observations and theories of stellar pulsations. This new development comes basically from the discovery of pulsations and oscillation-related phenomena in many stars, which were hitherto regarded as non-pulsating stars. They include our sun, white dwarfs, Ap stars, early-type O/B stars, and so on. The most important characteristics of oscillations in these stars are that their oscillations are usually multi-periodic with several modes of oscillations involved, while most of the classical pulsating variables, such as Cepheids and RR Lyrae stars, present only a single mode. The increasing number of pulsating modes in a star leads to the increase of the amount of information derived from the pulsation. Indeed, any single period of classical pulsating stars, which are thought to be pulsating with the radial fundamental mode or the first overtone, has provided us a measure of the mean density of the star, but in the case of the double-mode Cepheids, which are pulsating in two different modes, then the stellar mass can be determined as one more item of information from the pulsations in addition to the mean density. A more conspicuous case is our sun. The sun is oscillating in thousands of nonradial eigenmodes, and this richness in the number of identified modes leads to the great success of “helioseismology”, by which we can probe the invisible internal structure of the sun in detail. The discovery of very small amplitude pulsations, especially those found from observations of the sun as a star, has raised a question concerning the concept of pulsating and non-pulsating stars. We now expect that pulsations, at least with small amplitudes, will be ubiquitous among stars in general. This opens the possibility of a seismological approach to stars, and the research field probing the internal structure of stars in general is now called “asteroseismology”. We may not detect such large numbers of eigenmodes of individual distant stars even if they are oscillating, since we may not resolve the stellar image as in the case of the sun. However, we should be reminded that every one of the stars differs from all stars in its mass, and in its evolutionary stage. From the standpoint of the theory of stellar

oscillation, stars are like musical instruments that are able to oscillate in modes which differ from one star to another and change in delicate ways as a star evolves. Probing the invisible interior of distant stars in detail like in the case of the sun is thus worth doing. Asteroseismology is at the present moment still in its infancy, but it has the potential to develop into a major field of stellar physics.

2. Basics of stellar oscillations

In this section, I summarize the general properties of pulsations of stars. For more complete reviews on this subject, text books such as those by Ledoux and Walraven (1958); by Cox (1980); by Unno *et al.* (1989); and by Gough (1993) should be consulted. Recent progress has also been concisely described by Gautschi and Saio (1995, 1996).

2.1 Basic equations of linear oscillations

Stars are like musical instruments which have various modes of oscillations and tones. Such oscillations of a star are described by equations of hydrodynamics. We consider a spherically symmetric star as an unperturbed state upon which small perturbations of oscillation are superimposed. We assume that no magnetic field and no motions exist in the unperturbed state. We also assume in the linear theory that all perturbations are sufficiently small so that only terms in the first-order in perturbations are retained while those higher than the second are neglected. There are two different ways to express perturbations: the Eulerian form and the Lagrangian form. The Eulerian perturbation is defined as the perturbation of a physical quantity at a given position, denoted by prime, while the Lagrangian perturbation is defined by that for a given fluid element, denoted by a symbol δ . The connection between the Eulerian and Lagrangian perturbations f' and δf of any physical quantity f , respectively, is, to first order in the displacement of the fluid element,

$$\boldsymbol{\xi} \equiv \mathbf{r} - \mathbf{r}_0, \quad (1)$$

given by

$$\delta f = f' + \boldsymbol{\xi} \cdot \nabla f_0, \quad (2)$$

where $\mathbf{r}$ in equation (1) denotes the Lagrangian position variable of a given fluid element which is at $\mathbf{r} = \mathbf{r}_0$ in the equilibrium state, and the subscript 0 in equation (2) denotes the unperturbed, equilibrium state of f . (In what follows, we omit the subscript 0 for equilibrium quantities unless there is confusion.)

The set of linearized basic equations of hydrodynamics is derived as follows:

(a) Equation of continuity:

$$\frac{\partial \rho'}{\partial t} + \nabla \cdot (\rho \mathbf{v}') = 0. \quad (3)$$

(b) Equation of motion:

$$\frac{\partial \mathbf{v}'}{\partial t} = -\frac{1}{\rho} \nabla p' + \frac{\rho'}{\rho^2} \nabla p - \nabla \psi'. \quad (4)$$

(c) Poisson equation:

$$\nabla^2 \psi' = 4\pi G \rho'. \quad (5)$$

(d) Equation of thermal energy conservation:

$$\begin{aligned} \rho T \frac{\partial \delta S}{\partial t} &\equiv \rho T \frac{c_p}{v_T} \left\{ \frac{\partial (\delta \rho / \rho)}{\partial t} - \frac{1}{\Gamma_1} \frac{\partial (\delta p / p)}{\partial t} \right\} \\ &= \rho \varepsilon \left(\frac{\delta \rho}{\rho} + \frac{\delta \varepsilon}{\varepsilon} \right) - \delta(\nabla \cdot \mathbf{F}), \end{aligned} \quad (6)$$

where S denotes the specific entropy, ε the nuclear energy generation rate per unit mass, $\mathbf{F}$ the energy flux, $c_p \equiv T(\partial S / \partial T)_p$ the specific heat per unit mass at constant pressure, $v_T \equiv -(\partial \ln \rho / \partial \ln T)_p$, $\Gamma_1 \equiv (\partial \ln p / \partial \ln \rho)_S$, and the other symbols have their usual meanings (t : time, ρ : density, p : pressure, T : temperature, $\mathbf{v}$: velocity, ψ : gravitational potential, G : gravitational constant). Here we neglect viscosity since it is generally small in the stellar interior. The Lagrangian perturbation of the chemical composition is also neglected, since the periods of pulsations are much shorter than the time scales of nucleosynthesis and material diffusion. The Lagrangian variation of the velocity is the Stokes derivative of the displacement of the fluid element:

$$\delta \mathbf{v} = \left(\frac{\partial}{\partial t} + \mathbf{v} \cdot \nabla \right) \boldsymbol{\xi}. \quad (7)$$

Since we are assuming no velocity field in the unperturbed state, equations (2) and (7) lead to

$$\mathbf{v}' = \frac{\partial \boldsymbol{\xi}}{\partial t}. \quad (8)$$

2.2 Linear adiabatic wave equation: Eigenvalue problem

The temporal dependence of any perturbed quantities can be represented by $\exp(i\sigma t)$. The ratio of thermal to dynamical time scales is one of the most important parameters that govern the entropy perturbation of stellar pulsations. This fact is apparent in the dimensionless equation of energy conservation. By replacing $\partial/\partial t$ with $i\sigma$ after multiplying equation (6) by $4\pi r^3/L$, where L denotes the luminosity of the star and r , the distance from the stellar center, we obtain a dimensionless form of the linearized equation of thermal energy conservation,

$$i\omega \frac{\tau_{\text{th}}}{\tau_{\text{dyn}}} \frac{\delta S}{c_p} = \frac{4\pi r^3 \rho \varepsilon}{L} \left[\rho \varepsilon \left(\frac{\delta \rho}{\rho} + \frac{\delta \varepsilon}{\varepsilon} \right) - \delta(\nabla \cdot \mathbf{F}) \right]. \quad (9)$$

Here the thermal time scale τ_{th} and the dynamical time scale τ_{dyn} are defined as $\tau_{\text{th}} \equiv 4\pi r^3 \rho c_p T / L$, $\tau_{\text{dyn}} \equiv (R^3 / GM)^{1/2}$, respectively, where M and R denote the mass and the radius of the star, respectively, and ω denotes the dimensionless angular frequency normalized by τ_{dyn}^{-1} ; $\omega \equiv \sigma / (R^3 / GM)^{-1/2}$. Equation (9) indicates that for a given

energy excess and for a given c_p the entropy perturbation is smaller for larger $\tau_{\text{th}}/\tau_{\text{dyn}}$. The ratio $\tau_{\text{th}}/\tau_{\text{dyn}}$ is very large in the interior of a star except near the stellar surface. In most of the stellar interior, pulsations are then almost adiabatic, and the adiabatic pulsation is, in many cases, a very good approximation for the actual stellar pulsation. In this case, $\delta S = 0$, and hence

$$c^2 \frac{\partial \rho'}{\partial t} - \frac{\partial p'}{\partial t} = -\mathbf{v}' \cdot (c^2 \nabla \rho - \nabla p), \quad (10)$$

where $c \equiv (\Gamma_1 p / \rho)^{1/2}$ denotes the sound speed.

By operating $\partial/\partial t$ to the equation of motion (4), and eliminating $\partial \rho'/\partial t$ and $\partial p'/\partial t$ with the help of equations (3) and (10), we deduce the basic equation in terms of an operator $\mathcal{L}$ with respect to the displacement vector $\boldsymbol{\xi}$

$$\begin{aligned} \mathcal{L}(\boldsymbol{\xi}) \equiv & \frac{1}{\rho^2} \nabla p \nabla \cdot (\rho \boldsymbol{\xi}) - \frac{1}{\rho} \nabla (\boldsymbol{\xi} \cdot \nabla p) - \frac{1}{\rho} \nabla (c^2 \rho \nabla \cdot \boldsymbol{\xi}) \\ & + \nabla \left\{ G \int \frac{\nabla \cdot [\rho \boldsymbol{\xi}] d^3 \mathbf{x}}{|\mathbf{x} - \mathbf{r}|} \right\} = \sigma^2 \boldsymbol{\xi}. \end{aligned} \quad (11)$$

This equation, together with boundary conditions, forms an eigenvalue problem with the eigenvalue σ^2 .

The operator $\mathcal{L}$ is Hermitian in the case of vanishing pressure at the stellar surface (the zero-boundary condition):

$$\int_0^M \boldsymbol{\eta}^* \mathcal{L}(\boldsymbol{\xi}) dM_r = \int_0^M \boldsymbol{\xi}^* \mathcal{L}(\boldsymbol{\eta}) dM_r, \quad (12)$$

where M_r denotes the mass inside the radius r ; that is, $dM_r = 4\pi \rho r^2 dr$. Hence the eigenvalues, σ^2 , are real; the eigenfunctions form a normalized orthogonal set:

$$\int_0^M \boldsymbol{\xi}_q^* \cdot \boldsymbol{\xi}_{q'} dM_r = \delta_{qq'}, \quad (13)$$

where the subscripts attached to $\boldsymbol{\xi}$ label the eigenmode, and $\delta_{qq'}$ denotes the Kronecker delta.

The normal modes with non-zero eigenfrequencies σ in a spherically symmetric stars are characterized by the eigenfunctions that are proportional to the spherical harmonics, $Y_l^m(\theta, \phi)$ ($l = 0, 1, 2, \dots$; $m = 0, \pm 1, \dots, \pm l$), and the displacement eigenfunction is given by

$$\boldsymbol{\xi} = \left[\xi_r(r) \mathbf{e}_r + \xi_h(r) \left(\mathbf{e}_\theta \frac{\partial}{\partial \theta} + \mathbf{e}_\phi \frac{1}{\sin \theta} \frac{\partial}{\partial \phi} \right) \right] Y_l^m(\theta, \phi) \exp(i\sigma t). \quad (14)$$

The degree l represents the total number of nodal lines which divide the stellar surface into zones of opposite phase, while the azimuthal order m indicates the number of nodal lines which are longitudinal. The normal modes belonging to a spherical harmonic are further distinguished by the number of nodes, n , in the radial component of displacement from the center to the surface of a star. The normal modes are then labelled by the three quantum numbers: the radial order n , the degree l , and the

azimuthal order m . The eigenfrequencies, however, depend on the radial order n and the degree l but are degenerated by $(2l + 1)$ -folds in the azimuthal order m in the case of a non-rotating, non-magnetic spherical star. This degeneracy is lifted if rotation or the magnetic field of the star is taken into account.

There exists a trivial solution to the basic equation (11) of oscillations besides the spheroidal modes given by equation (14). In this solution, the eigenfrequency σ and all scalar variables are zero, but the displacement vector is non-zero and given by

$$\xi = -\frac{1}{\sqrt{l(l+1)}} \left(-\mathbf{e}_\theta \frac{1}{\sin \theta} \frac{\partial}{\partial \phi} + \mathbf{e}_\phi \frac{\partial}{\partial \theta} \right) Y_l^m(\theta, \phi). \quad (15)$$

These toroidal modes are steady eddy motions and not of much interest in the case of a non-rotating, non-magnetic spherical star, but they become oscillatory modes in the case of a rotating and/or magnetic star.

It is known that the set of eigenfunctions which consist of the spheroidal modes and the toroidal modes is complete (Eisenfeld 1969), and hence any oscillation can be expressed in terms of a series expansion of this set.

Taking the scalar product of equation (11) with $\rho \xi^*$ and integrating it throughout the stellar volume, we obtain an expression of the eigenvalue in terms of the eigenfunction:

$$\sigma^2 = \int_0^M \xi^* \cdot \mathcal{L}(\xi) dM_r \quad (16)$$

From this expression, it can be shown that the eigenfrequencies obey a variational principle (Chandrasekhar 1964):

$$\delta \sigma^2 = \int_0^M \xi^* \cdot \delta \mathcal{L}(\xi) dM_r \quad (17)$$

The variational principle indicates that eigenfrequencies are determined more accurately than the eigenfunctions, and this fact was utilized to improve eigenfrequencies before the modern computer age (Rayleigh–Ritz method). The variational principle recently came to light again as a basis of the inverse problem of the oscillations, in which the stellar structure is determined from the eigenfrequencies. In this case, $\delta \sigma^2$ in the left-hand-side of equation (17) is regarded as the difference between the observed frequency and the theoretical frequency of the same mode of a model of the star, and is a known quantity. On the other hand, the right-hand-side is dependent on the difference between the equilibrium quantities of the true star and those of the model, which is an unknown to be solved. From equation (17) for many observed modes in the case of the sun, we have well determined the sound speed profile and the density profile in the solar interior, from which we deduce other physical parameters in the solar interior.

2.3 Linear adiabatic radial pulsation

In the case of $l = 0$ (radial pulsation), the displacement vector has only the radial component

$$\xi = \xi(r) \mathbf{e}_r \exp(i\sigma t), \quad (18)$$

and equation (11) is reduced to a Sturm-Liouville type equation:

$$\frac{d}{dr} \left(\Gamma_1 p r^4 \frac{d\zeta}{dr} \right) + r^3 \frac{d}{dr} [(3\Gamma_1 - 4)p] \zeta = \sigma^2 \rho r^4 \zeta, \quad (19)$$

where $\zeta \equiv \xi/r$. Hence, the eigenfrequencies are lower bounded:

$$\sigma_1^2 < \sigma_2^2 < \sigma_3^2 < \dots \quad (20)$$

The restoring force is the pressure force, and then the pulsation modes are acoustic waves.

2.4 Linear adiabatic nonradial oscillations

For non-radial pulsations ($l \neq 0$), a star pulsates in such a way as to deviate from a spherical shape. Two different kinds of restoring forces (the pressure force and the buoyancy force) operate, and there exist, therefore, two different kinds of modes: acoustic modes (p -modes) and gravity modes (g -modes). To discuss general properties of non-radial oscillations the Cowling (1941) approximation, in which the Eulerian perturbation to the gravitational potential is neglected ($\psi' = 0$), is appropriate to use, since it has a sufficient degree of accuracy and simplifies the treatment greatly. The accuracy of this approximation is quite good for modes with large values of n and l . Using the Cowling approximation allows us to ignore the last term in the middle of equation (11). Note that this equation is not of the Sturm-Liouville type.

It is instructive to consider the wave propagation in a plane isothermal atmosphere under a constant gravitational field g , in which the pressure p and density ρ decrease with height z as $\propto \exp(-z/H)$. The perturbation is described by

$$\xi \propto \exp\left(\frac{z}{2H}\right) \exp[i(\sigma t + k_h x + k_z z)], \quad (21)$$

where x stands for the horizontal coordinates. The exponentially growing factor with height in equation (21) arises so as to conserve wave energy in the vertical direction since the density ρ in atmosphere decreases with height. The angular frequency σ and the horizontal and vertical wavenumbers k_h and k_z must then satisfy a dispersion relation which is given by

$$\sigma^4 - \sigma^2[(k_h^2 + k_z^2)c^2 + \sigma_{ac}^2] + k_h^2 c^2 N^2 = 0, \quad (22)$$

where $N \equiv [gd \ln(p^{1/\Gamma_1}/\rho)/dz]^{1/2}$ is the Brunt-Väisälä frequency and $\sigma_{ac} \equiv c/(2H)$ is the acoustic cut-off frequency. If σ and k_h are given, this dispersion relation determines k_z^2 . If $k_z^2 > 0$, waves can propagate vertically. On the other hand, if $k_z^2 < 0$, the energy density perturbation decreases exponentially with height if no wave flux is coming from above. The situation can be seen if we plot the relation (22) in the (k_h, σ) -diagram, which is divided into three regions (see Fig. 1): (i) $k_z^2 > 0$, where acoustic waves can propagate (region 'P' of Fig. 1), (ii) $k_z^2 > 0$, where gravity waves can propagate (region 'G' of Fig. 1), and (iii) $k_z^2 < 0$, where waves are evanescent.

Actual stars are by no means isothermal or plane under constant gravity. However, the diagnostic diagram is still helpful in approximately representing the wave nature

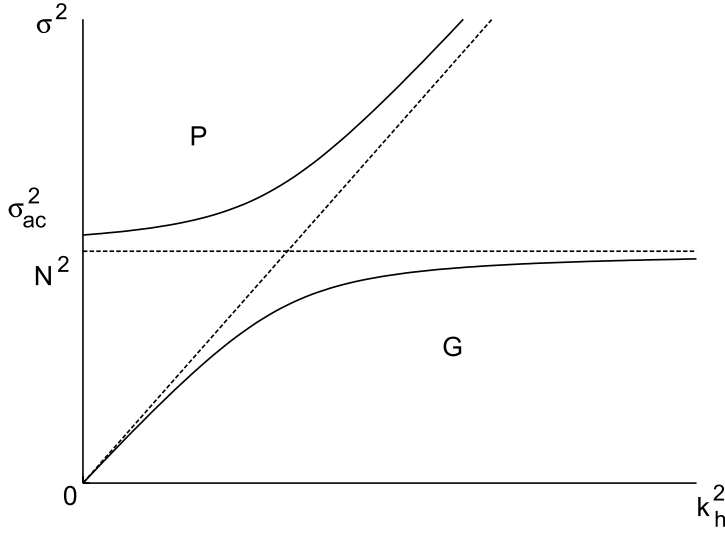


Figure 1. Diagnostic diagram for the plane isothermal atmosphere. The quantity k_z^2 is positive in the region 'P' and in the region 'G', while it is negative in the other regions.

of non-radial pulsations of stars. Let us substitute the form of solution (14) into the wave equation (11) under the Cowling approximation, and write the wave equation (11) symbolically

$$\frac{d}{dr} \begin{pmatrix} \xi_r \\ \xi_h \end{pmatrix} = \begin{pmatrix} \mathcal{A} & \mathcal{B} \\ \mathcal{C} & \mathcal{D} \end{pmatrix} \begin{pmatrix} \xi_r \\ \xi_h \end{pmatrix}. \quad (23)$$

With the transformation of variables ξ_r and ξ_h to new variables u and v defined by $u \equiv \mathcal{B}^{-1/2} \xi_r$ and $v \equiv \mathcal{C}^{-1/2} \xi_h$, the wave equation results in

$$\frac{d^2 u}{dr^2} + \frac{1}{c^2 \sigma^2} [\sigma^2 - \Phi_p(r)] [\sigma^2 - \Phi_g(r)] u = 0, \quad (24)$$

where

$$\Phi_p \simeq \frac{l(l+1)c^2}{r^2} + \sigma_{ac}^2 \quad (25)$$

and

$$\Phi_g \simeq N^2 \quad (26)$$

are functions of r . If we assume $u \propto \exp(\int k_r dr)$, the wave propagates radially as an acoustic wave if $\Phi_g, \Phi_p < \sigma^2$, and as a gravity wave if $\sigma^2 < \Phi_g, \Phi_p$. The wave is evanescent if $\Phi_g < \sigma^2 < \Phi_p$ or $\Phi_p < \sigma^2 < \Phi_g$. Note that equation (24) tends to a Sturm–Liouville type as $\sigma^2 \rightarrow \infty$ or $\sigma^2 \rightarrow 0$. The angular eigenfrequency σ and the radial dependence of the eigenfunction $u(r)$ and $v(r)$ are described by the radial order n and the degree l of the mode, but are independent of the azimuthal order m for the case of a spherical star in its unperturbed state. For a given l , the azimuthal

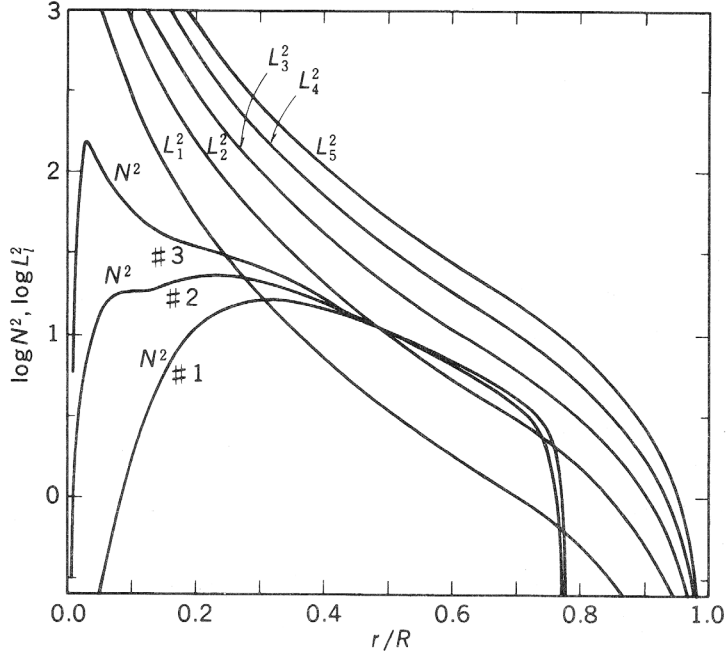


Figure 2. Propagation diagram for models of a $1 M_{\odot}$ star at the ZAMS stage (1) and at the advanced evolutionary stages (2 and 3). Model 2 has an internal structure close to that of the present sun. The quantities N^2 and $L_l^2 \equiv l(l+1)c^2/r^2$ are measured in units of GM/R^3 . The L_l^2 -curves are of model 2, but they do not differ much for different models.

order m takes integer values from $-l$ to l , and thus the nonradial modes are $(2l+1)$ -fold degenerate. This degeneracy is lifted if rotation or the magnetic field of the star is taken into account.

The wave propagation feature is easily seen in the propagation diagram (Fig. 2), on which $\Phi_p(r)$ and $\Phi_g(r)$ are plotted as functions of r . Let us consider the propagation diagrams for the ZAMS model and later core-hydrogen-burning phase models of a star. Note that the squared Brunt-Väisälä frequency is negative in the convective zone and the peak frequency increases with evolution in the μ -gradient zone located just outside the nuclear reacting core, since N^2 includes a term of the gradient of the mean molecular weight $\nabla_{\mu} \equiv d \ln \mu / d \ln p$, which increases with evolution:

$$N^2 = g^2 \frac{\rho}{p} \left[v_T (\nabla_{\text{ad}} - \nabla) + \left(\frac{\partial \ln \rho}{\partial \ln \mu} \right)_{p,T} \nabla_{\mu} \right], \quad (27)$$

where $g \equiv GM_r/r^2$, $\nabla_{\text{ad}} \equiv (\partial \ln T / \partial \ln p)_S$ and $\nabla \equiv d \ln T / d \ln p$. The μ -gradient zone acts like a potential well that traps gravity waves. In the case of a star of $M \geq 2 M_{\odot}$, with the increase of N^2 in the μ -gradient zone, there appears an evanescent zone, for modes in the frequency range between the two maxima of the N^2 -curve, and these modes possess the dual character such that they behave like gravity waves in the μ -gradient zone, but like acoustic waves in the outer part of the star. The evolutionary development of the μ -gradient zone causes the star to have a structure similar to two

potential wells separated by an evanescent zone. One of them is a potential well for a gravity-wave oscillation trapped in the interior and the other is a potential well for an acoustic-wave oscillation trapped in the envelope. The dichotomy is particularly distinguishable in the case of high degree l , but it is obscure for low degree modes such as $l = 1$ or 2 . Hence, though there is no chance to detect the high degree modes well trapped in the μ -gradient zone, it may be possible to detect the low degree modes oscillating even in the μ -gradient zone. Detection of such modes will provide us with information about the evolutionary stage of the star and the structure of the deep interior of the star.

When a pair of the frequencies, one for the mode trapped in the interior and the other for the mode trapped in the envelope, approach each other, these modes interact with each other, and they exchange the oscillation properties without equalizing their frequencies. If we plot the evolutionary change of these frequencies as a function of evolutionary time, these frequencies approach once each other and then separate without crossing. This phenomenon is called “avoided crossing” in the community of stellar pulsation, while a similar phenomenon is called “nearly kissing” in the community of geophysics.

After the central hydrogen is exhausted and the core becomes almost isothermal, the squared Brunt–Väisälä frequency still becomes higher near the center, while farther out it becomes negative since the envelope becomes convectively unstable. As the star evolves further into the white dwarf stage, the core becomes degenerate. The Brunt–Väisälä frequency becomes zero in the completely degenerate and chemically homogeneous core, and N^2 is positive only near the stellar surface in the case of white dwarfs. In this case, the gravity modes are the oscillations trapped near the surface while the acoustic waves can penetrate into the deep interior, contrary to the case of non-degenerate stars. As the white dwarf cools and the luminosity becomes lower, the frequencies of gravity modes become lower.

3. Excitation mechanisms of stellar pulsations: Thermal overstability

As discussed in the previous section, adiabatic oscillations in a non-rotating and non-magnetic spherical star are strictly periodic. Oscillations in nature are, however, inevitably nonadiabatic. This corresponds to the fact that the eigenfrequencies and eigenfunctions are no longer purely real in the mathematical description of linear nonadiabatic oscillations. In order for a particular mode to be excited, some excitation mechanism must be present in a region where the amplitudes of eigenfunctions are large. The pulsational stability is a problem of the exchange from one type of energy to another in a system. Here we consider the linear stability problem to investigate the restoring forces and excitation mechanisms of oscillation. For more complete reviews on this subject, the text book by Unno *et al.* (1989) should be consulted.

3.1 Energy equation and the work integral

We arrange the basic equations (3), (4), and (6) in such a way as to give the partial time derivatives $\partial \mathbf{v}' / \partial t$, $\partial p' / \partial t$, and $\partial (p' - c^2 \rho') / \partial t$; multiply these equations by $\rho \mathbf{v}'$, $p' / (\Gamma_1 p)$, and $g^2 (p' - c^2 \rho') / (\rho c^4 N^2)$, respectively; and add them. After somewhat lengthy manipulations (Unno *et al.* 1989), the result turns out to be

$$\frac{\partial}{\partial t}(\rho e_W) + \nabla \cdot \mathbf{F}_W = \rho \delta T \frac{\partial \delta S}{\partial t} - \Phi' \frac{\partial \rho'}{\partial t}, \quad (28)$$

where wave energy per unit mass, e_W , is defined as

$$e_W = \frac{1}{2} \left\{ \mathbf{v}'^2 + \left(\frac{p'}{\rho c} \right)^2 + \left(\frac{g}{N} \right)^2 \left[\left(\frac{p'}{\Gamma_1 p} - \frac{\rho'}{\rho} \right)^2 - \frac{\nabla}{\nabla_{\text{ad}}} \left(v_T \frac{\delta S}{c_p} \right)^2 \right] \right\}, \quad (29)$$

and

$$\mathbf{F}_W = p' \mathbf{v}' + \rho \mathbf{v}' \Phi'. \quad (30)$$

We can replace $\partial/\partial t$ by the Stokes derivative d/dt in equation (28) because the difference appears only in higher order terms. The second term in the right-hand side of equation (28) can be incorporated into the global wave energy E_W by integrating equation (28) over the whole volume of the star; *i.e.*,

$$\frac{dE_W}{dt} = \int_0^M \delta T \frac{d\delta S}{dt} dM_r - \int_S \mathbf{F}_W \cdot d\mathbf{S}, \quad (31)$$

where

$$E_W = \int_0^M \left(e_W + \frac{1}{2} \frac{\rho'}{\rho} \Phi' \right) dM_r. \quad (32)$$

This equation states that change in the wave energy of a star is caused by non-adiabatic processes in the interior and by the outgoing wave flux $\mathbf{F}_W$ at the surface. The first term in e_W is the kinetic energy and the other terms represent the potential energies corresponding to the various restoring forces. If there is a negative potential energy, monotonically unstable modes could exist, because the kinetic energy can increase without changing the total energy. For example, a dynamically unstable convective mode arises if $N^2 < 0$ somewhere in the stellar interior, due to the exchange of energy between the first and the third term in the right-hand side of equation (29). In the case of an acoustic wave, the second term in e_W is dominant, and the wave energy averaged over one period is equi-partitioned into the kinetic energy $(1/2)\mathbf{v}'^2$ and the acoustic potential energy $(1/2)(p'/\rho c)^2$. When the third term is dominant, the internal gravity wave ($N^2 > 0$) or the convective mode ($N^2 < 0$) appears. The term proportional to $(\delta S)^2$ is related to nonadiabatic effect. Although this term cannot be discussed independently of the right-hand side of equation (28), equation (29) suggests that if this term is so large as to make the square bracket negative, the convective mode ($N^2 < 0$) could become oscillatory.

Integrating equation (31) over one period of oscillation, we obtain

$$\begin{aligned} W &\equiv \oint \frac{dE_W}{dt} dt = \oint dt \left[\int_0^M \delta T \frac{d\delta S}{dt} dM_r - \int_S \mathbf{F}_W \cdot d\mathbf{S} \right] \\ &= \oint dt \left[\int_0^M \frac{\delta T}{T} \delta \left(\varepsilon - \frac{1}{\rho} \nabla \cdot \mathbf{F} \right) dM_r \right] \\ &\quad - \oint dt \left[\int_{r=R} (p' + \rho \Phi') \mathbf{v}' \cdot d\mathbf{S} \right], \end{aligned} \quad (33)$$

where $\oint$ indicates the integration over one period of oscillation. This quantity is defined as the work integral W and it means the increase of the global wave energy over one period of oscillation. In the case of the zero-boundary condition, that is, if the pressure goes to zero at the surface, the work integral W turns to be the first term in the second line of equation (33) (Eddington 1926). If W is positive, the star is vibrationally unstable. The energy increase is provided from photon energy, which is originally produced by nuclear reactions in the core.

Equation (33) indicates that if a finite value can artificially be given to the surface pressure, a strictly periodic oscillation can be produced by equating the first term in the right-hand side to the second term. This corresponds to a strictly periodic imaginary oscillation, in which all the excitation effect is consumed by the work done on the artificially placed matter above the stellar surface. The work integral W can also be defined as the amount of energy which must be removed from the star in order for the star to oscillate strictly periodically.

The work integral W is related to the growth rate $\sigma_{\Im}$ of amplitude of oscillation, which is the imaginary part of σ ($\sigma = \sigma_{\Re} + i\sigma_{\Im}$):

$$\sigma_{\Im} = -\frac{1}{2} \frac{W/E_W}{\Pi}, \quad (34)$$

where Π denotes the period. The factor $1/2$ comes from the fact that the energy is proportional to the square of amplitude. We consider that $|\sigma_{\Im}/\sigma_{\Re}| \ll 1$, since the ratio is of the order of magnitude of the dynamical to thermal time-scale ratio. Then, $\Pi = 2\pi\sigma_{\Re}^{-1}$. The total energy of oscillation E_W is twice the time average of the kinetic energy since there is equi-partition of the kinetic and potential energies of oscillation in the time average:

$$E_W = \frac{\sigma_{\Re}^2}{2} \int_0^M |\xi|^2 dM_r. \quad (35)$$

Then, we have

$$\sigma_{\Im} = -\frac{1}{2\pi\sigma_{\Re}} W \left[\int_0^M |\xi|^2 dM_r \right]^{-1}. \quad (36)$$

3.2 κ -mechanism

In quasi-adiabatic analysis the work integral W is estimated by using adiabatic eigenfunctions as well as adiabatic relations. In this case we should evaluate the work integral W by using the expression given in equation (33). For the sake of simplicity, in what follows we neglect convection and the surface energy loss. Let us decompose W into W_N and W_F , which are related to the perturbations of nuclear energy generation rate and radiative flux, respectively:

$$W = W_N + W_F \quad (37)$$

with

$$W_N = \frac{\pi}{\sigma_{\Re}} \int_0^M \frac{\delta T^*}{T} \delta \varepsilon dM_r \quad (38)$$

and

$$W_F = \frac{\pi}{\sigma_{\mathfrak{R}}} \int_0^M \frac{\delta T^*}{T} \left[-\frac{d\delta L_r}{dM_r} + \frac{l(l+1)}{d \ln T / d \ln r} \frac{F}{\rho r} \frac{\delta T}{T} \right. \\ \left. + l(l+1) \left(\frac{\xi_h}{r} \frac{dL_r}{dM_r} - \frac{\xi_r}{r} \frac{F}{\rho r} \right) \right] dM_r, \quad (39)$$

where L_r and $F \equiv L_r/(4\pi r^2)$ denote the radiative luminosity at the radius r and the radiative flux, respectively. Here we neglected the surface integral $\int \mathbf{F}_W \cdot d\mathbf{S}$ representing the energy loss by the outgoing wave which vanishes for the trapped oscillation.

The contribution of the first term in the square bracket in the right-hand-side of equation (39) is positive if a positive value of δT is associated with a decrease of δL_r towards the surface. This means that instability may occur if matter gains energy at the compression phase and loses energy during expansion. It is instructive to modify further the expression for W_F to a form in which all the terms are proportional to $(\delta T/T)^2$ or $[d(\delta T/T)/dr]^2$. After some manipulations (Unno *et al.* 1989), we obtain an approximate expression for W_F :

$$\frac{\sigma_{\mathfrak{R}}}{\pi} W_F \simeq -\frac{1}{2} \left[\alpha_1 L_r \left(\frac{\delta T}{T} \right)^2 \right]_{r=R} - \frac{1}{2} \int_0^R dr \left(\frac{\delta T}{T} \right)^2 \frac{d}{dr} (\alpha_1 L_r) \\ + \int_0^R dr L_r \frac{H_p}{\nabla_{\text{ad}}} \left(\frac{\nabla - \nabla_{\text{ad}}}{\nabla} - \frac{c_1 \omega^2 - U}{\alpha_0} \right) \\ \times \left\{ \left[\frac{d}{dr} \left(\frac{\delta T}{T} \right) \right]^2 + \frac{l(l+1)}{r^2} \left(\frac{\delta T}{T} \right)^2 \right\} \\ + l(l+1) \int_0^R dr \left(\frac{\delta T}{T} \right)^2 \frac{1}{\nabla_{\text{ad}}} \left[\frac{4-V}{\alpha_0 V} \left(\frac{1}{c_1 \omega^2} \frac{dL_r}{dr} - \frac{L_r}{r} \right) \right. \\ \left. + \frac{c_1 \omega^2 - U}{V \alpha_0 c_1 \omega^2} \frac{dL_r}{dr} + \frac{1}{2} \frac{d}{dr} \left(\frac{c_1 \omega^2 L_r + dL_r/d \ln r}{c_1 \omega^2 \alpha_0 V} \right) \right]. \quad (40)$$

Here α_0 and α_1 are defined as

$$\alpha_0 = 4 - U - \frac{l(l+1)}{c_1 \omega^2} + c_1 \omega^2 \quad (41)$$

and

$$\alpha_1 \equiv 4 - \frac{1}{\nabla_{\text{ad}}} - \kappa_T - \frac{\kappa_\rho}{\Gamma_3 - 1} + \left(1 - \frac{l(l+1)c^2}{\Gamma_1 \sigma_R^2 r^2} \right) \left(\frac{c_1 \omega^2 - U}{\alpha_0 \nabla_{\text{ad}}} \right), \quad (42)$$

respectively, where

$$\begin{aligned}
 c_1 &\equiv \left(\frac{r}{R}\right)^3 / \left(\frac{M_r}{M}\right), \\
 \omega^2 &\equiv \frac{\sigma_{\mathfrak{R}}^2}{(GM/R^3)}, \\
 U &\equiv \frac{d \ln M_r}{d \ln r}, \\
 V &\equiv -\frac{d \ln p}{d \ln r}, \\
 H_p &\equiv -\frac{dr}{d \ln p}, \\
 \Gamma_3 - 1 &\equiv \left(\frac{\partial \ln T}{\partial \ln \rho}\right)_s, \\
 \kappa_T &\equiv \left(\frac{\partial \ln \kappa}{\partial \ln T}\right)_\rho, \\
 \kappa_\rho &\equiv \left(\frac{\partial \ln \kappa}{\partial \ln \rho}\right)_T,
 \end{aligned}$$

and where κ denotes the opacity.

The first and the second terms in the right-hand side of equation (40) describe the κ -mechanism (Baker and Kippenhahn 1962). In the outer envelope in the radiative equilibrium, L_r is constant and the κ -mechanism works for driving an oscillation if

$$\frac{d}{dr} \left(\kappa_T + \frac{\kappa_\rho}{\Gamma_3 - 1} \right) > 0. \quad (43)$$

If a region in the stellar envelope satisfies this condition, radiative flux from the stellar interior is blocked by the effect of the temperature and density dependence of opacity. The blocked energy is converted to the energy of the oscillation. The value of κ_T increases in the inner part of an ionization zone and decreases in the outer part. Therefore, the excitation and damping zones due to the κ -mechanism are located in, respectively, the inner and outer parts of the ionization zone. In an ionization zone an adiabatic exponent ($\Gamma_3 - 1$) (> 0) is minimum. This spatial variation of ($\Gamma_3 - 1$) enhances the effect of the κ -mechanism. This effect is sometimes called the γ -mechanism (e.g., Cox *et al.* 1966).

It should be noted here that we would expect that the strong damping of the outer portions of an ionization zone would largely cancel the strong driving of the inner portions. However, because of non-adiabatic effects, the non-adiabatic eigenfunctions can be quite different from the adiabatic ones in layers sufficiently far out in a star. It is this lack of equality between these two quantities that leads to the overall effect of the ionization zone on the stability of the star. The strongly non-adiabatic exterior is

separated from the quasi-adiabatic interior by a transition region of which location in the stellar envelope is determined by

$$\tau_{\text{th}}/\tau_{\text{dyn}} \sim 1, \quad (44)$$

where τ_{th} and τ_{dyn} are defined in section 2.2. Since the decrease of τ_{th} with r is very rapid, the location of the transition region depends on the oscillation modes only weakly. Depending on the equilibrium models and the oscillation mode, the transition region can be located in the hydrogen ionization zone or in the helium ionization zone or in the ionization of carbon and oxygen.

It is well known after Zhevakin (1953) that the κ -mechanism in the hydrogen and helium ionization zones is responsible for the Cepheid instability strip of radial pulsations which extends from the classical Cepheids in the giant region to δ Scuti variables near the main sequence. Moreover, the κ -mechanism in the hydrogen ionization zone has been investigated as a possible excitation mechanism for the observed nonradial g -modes of the variable DA white dwarfs, namely DAV stars. Also, overstable g -modes of variable DB white dwarfs (helium envelope; DBV stars) are thought to be excited by the κ -mechanism in the second helium ionization zone. It should be noted, however, that an alternative convective mechanism has been proposed and the excitation mechanism for DAV and DBV stars is still controversial (Brickhill 1991; Goldreich & Wu 1999). The κ -mechanism in the hydrogen ionization zone in the atmosphere of Ap stars has been extensively investigated by Balmforth *et al.* (2001) and Cunha (2002) as a possible excitation mechanism for rapid oscillations in roAp stars [see also Saio (2005)].

The new opacity data which were recalculated for the OPAL project (Rogers & Iglesias 1992; Iglesias & Rogers 1993, 1996) and for the OP project (Seaton *et al.* 1994) show a significant enhancement over the old Los Alamos opacity library in the temperature range of $T \sim 100,000\text{--}300,000$ K. This feature called the “Z-bump” is caused by intra-M-shell transitions mainly in Fe and by fine-structure transitions in Mg, Cr, and Ni. Based on the new opacity data, one could eventually, after many years of futile attempts, identify driving mechanisms for pulsations in early-type stars, that is, β Cephei stars and slowly pulsating B-type (SPB) stars (Kiriakidis *et al.* 1992; Moskalik & Dziembowski 1992; Gautschy & Saio 1993; Dziembowski *et al.* 1993). It was also predicted that the κ -mechanism associated with a local enrichment of iron caused by diffusion in hot subdwarf B (sdB) stars works to drive pulsations in these stars. Independent discovery of real sdB pulsators was sensational (see review, Charpinet *et al.* 2001).

The κ -mechanism associated with the ionization of carbon and oxygen is also possible. In the case of very hot pre-white-dwarf stars, including planetary nebula nuclei, of which hydrogen-rich matter has been lost, carbon/oxygen matter produced by helium burning could be located in the temperature region of $T \sim 10^6$ K. In such a case the κ -mechanism associated with the ionization of the K-shell electrons of carbon/oxygen can excite global oscillations. DOV stars showing nonradial pulsations are interpreted as such stars.

3.3 δ -mechanism

The second line of equation (40) arises from the radiative diffusions of the thermal energy of oscillating gaseous elements. For high order g -modes $(c_1\omega^2 - U)/\alpha_0$ is very

small. This indicates a superadiabatic region ($\nabla > \nabla_{\text{ad}}$) excites high order g -modes. This excitation effect is called the δ -mechanism in the text book by Unno *et al.* (1989). This excitation effect was first demonstrated by Cowling (1957), who showed the overstability of g -modes for the superadiabatic stratification of plasma stabilized by the existence of horizontal magnetic fields. If the superadiabatic region is stabilized by a spatial gradient of the mean molecular weight, this mechanism corresponds to Kato's (1966) mechanism. On the other hand, this mechanism does not exist for p -modes or radial pulsations, because in such cases $[(\nabla - \nabla_{\text{ad}})\nabla^{-1} - (c_1\omega^2 - U)\alpha_0^{-1}]$ reduces to $\sim -\nabla_{\text{ad}}/\nabla$. This term causes radiative damping for p -modes and radial pulsations irrespective of the sign of $(\nabla - \nabla_{\text{ad}})$.

A better understanding of this mechanism may be obtained by a local analysis. To make the analysis simple we use the Boussinesq approximation, in which density is treated as a constant in the equation of continuity and in the equation of motion except the buoyancy term. Let us suppose a plane-parallel, gravitationally stratified layer of fluid in hydrostatic and radiative equilibrium. We consider the higher overtone of gravity waves whose wavelengths in the vertical direction are much shorter than the density and pressure scale heights. This implies that the density variations are very small, and this is the situation appropriate for the Boussinesq approximation, in which the density variations are taken into account only in the term of the buoyancy, and the pressure perturbation p' is neglected in the energy equation. If the temporal and spatial dependence of the perturbation of an arbitrary quantity f is taken as

$$f' \propto \exp[i(k_h x + k_z z) + st], \quad (45)$$

where the gravity is the z -direction, then, the basic equations of oscillations lead to the dispersion relation

$$s^3 + \frac{Kk^2}{c_p \rho} s^2 + \frac{k_h^2}{k^2} N^2 s + \frac{k_h^2}{k^2} \frac{g}{H_p} \nabla_\mu \frac{Kk^2}{c_p \rho} = 0, \quad (46)$$

where $k^2 \equiv k_h^2 + k_z^2$, $K \equiv (4acT^3)/(3\kappa\rho)$ is the radiative conductivity, and g denotes the gravitational acceleration. The necessary and sufficient conditions for stability is $\Re(s) < 0$. Disregarding a root for the secular mode, we find two roots in the first approximation of small non-adiabaticity

$$s_{1,2} \simeq \frac{Kk^2}{2c_p p} v_T g^2 \frac{\nabla - \nabla_{\text{ad}}}{N^2} \pm i \frac{k_h}{k} N, \quad (47)$$

where N means the Brunt-Väisälä frequency given by equation (27). The dynamical instability occurs if $N^2 < 0$, that is, if

$$v_T (\nabla - \nabla_{\text{ad}}) - \nabla_\mu > 0, \quad (48)$$

which is often called the Ledoux (1947) criterion for the convective instability. In the dynamically stable case, the medium is stable against the ordinary convection since $N^2 > 0$, but perturbation grows oscillatory with increasing amplitude if

$$\nabla - \nabla_{\text{ad}} > 0 \quad (49)$$

because the real part of s in equation (47) is positive. The physical cause of this overstability is that the radiative heat exchange brings about an asymmetry in the

oscillatory motion in such a way that an oscillatory element overshoots its equilibrium position with an increasing velocity.

The favorable condition for this kind of overstability is realized in the μ -gradient zone around the convective core in massive stars, and it has been confirmed by a global stability analysis that some g -modes trapped in the μ -gradient zone are in fact overstable due to this mechanism (Shibahashi and Osaki 1976). In the presence of the uniform magnetic field, the condition for the dynamical stability is modified so that $k_h^2 N^2 / k^2 + (\mathbf{B} \cdot \mathbf{k})^2 / (4\pi\rho) > 0$, but the overstability condition (49) is unchanged. Shibahashi (1983) proposed this kind of overstability due to the superadiabatic layer in the magnetic stars' envelope as a possible excitation mechanism of roAp stars.

3.4 DB gap and pulsation of white dwarfs

In terms of surface composition, white dwarfs come in two major subgroups: those with essentially pure hydrogen atmospheres (DAs), which constitute about 80% of all white dwarfs, and those with essentially pure helium atmospheres (DBs). The ratio of DA to DB white dwarfs is temperature sensitive. There seem no DBs between 45,000 K and 30,000 K, and this exclusion zone is known as the "DB gap" (Fontaine and Wesemael 1991). Since the temperatures of the upper and the lower bounds of the DB gap coincide with the effective temperatures where the He II/III and He I/II convection zones show up, respectively, convective mixing is suspected of the cause of appearance of DBs outside the DB gap. Inversely, chemical separation due to gravitational settling in the convectively stable atmosphere is suspected of the cause of the presence of the DB gap. If one adopts this scenario, the DBs near the boundaries of the DB gap are expected to have a super-adiabatic layer which is nonetheless convectively stable due to the μ -gradient caused by chemical separation. Such DBs may be pulsationally unstable and then a new type of DB variables is expected (Shibahashi, in preparation).

4. Stochastic excitation of solar-like oscillations

Mode stability is governed not only by the perturbations in the radiative fluxes but also by the perturbations in the turbulent convective fluxes, which have been completely ignored in the previous section for the sake of simplicity. This is particularly true for the sun and solar-like stars, in which energy is mainly transported by convection in their outer envelopes. The principle source of excitations of the observed oscillations in the sun was controversial for many years. If solar p -modes were thermally overstable, some nonlinear mechanism must limit their amplitudes to the low values that are observed. But no such mechanism has been found. Alternatively, it is now widely believed that the observed low-amplitude oscillations in the sun are determined by the local balance between the stochastic driving due to the acoustic radiation by turbulent convection and the damping of the intrinsically stable p -modes. According to this view, the turbulent convection generates acoustic noise, and acoustic noise in the sun's resonant cavity results in the excitation of the cavity's normal modes. The kinetic energy of modes is stochastically supplied by the turbulence, and the radiation works to damp the modes. This view is extended to solar-like stars in general, and then detection of solar-like oscillations in late-type stars are vigorously attempted. In fact, detection of

solar-like oscillations in these stars has been accomplished in the last few years, after decades of disappointing attempts (see Kurtz 2004). Probing the invisible interior of stars is a dream come true. In this section, following mainly Osaki (1990), I first outline analytically the stochastic excitation mechanism. For more details, papers such as those by Goldreich and Keeley (1977); by Osaki (1990); by Balmforth (1992); by Goldreich *et al.* (1994); by Houdek *et al.* (1999); by Samadi and Goupil (2001); and by Chaplin *et al.* (2005) should be consulted. Then I briefly discuss the recent results based on numerical simulations.

4.1 Damping rate of oscillations and acoustic noise generation

Let us write the energy of a particular eigenmode, its angular frequency, and its energy dissipation rate as E_q , σ_q , and Γ_q , where the subscript, q , signifies a particular eigenmode. The energy of the mode is then given by

$$E_q = 2 \times \left\langle \frac{1}{2} \int_0^M v_q^2 dM_r \right\rangle = \left\langle \int_0^M v_q^2 dM_r \right\rangle, \quad (50)$$

where v_q denotes the velocity amplitude of the oscillation mode and the integration is performed over the whole mass of the star. Here the angular bracket means time average and we note that the energy of oscillation is the sum of the kinetic energy plus the potential energy. The equation of energy of the mode is then given by

$$\frac{dE_q}{dt} = G - \Gamma E_q, \quad (51)$$

where G denotes the acoustic power by turbulent convection available to the particular mode in our interest.

The damping rate of oscillation may be divided into two parts:

$$\Gamma = \Gamma_{\text{therm}} + \Gamma_{\text{dyn}}, \quad (52)$$

where Γ_{therm} and Γ_{dyn} stand for damping rates of oscillation that are caused through the thermal equation and the momentum equation, respectively. Furthermore, each of them may be written by several contributions by different causes. In fact, the thermal damping due to the non-adiabatic effects of oscillation may be divided into the damping due to radiative flux variation and that due to convective flux variation, and then Γ_{therm} may be written into two parts:

$$\Gamma_{\text{therm}} = \Gamma_{\text{rad}} + \Gamma_{\text{conv}}, \quad (53)$$

where Γ_{rad} and Γ_{conv} stand for the damping rates of oscillation due to radiative and convective flux variations. The thermal damping rate Γ_{therm} can be negative (*i.e.*, “negative” dissipation) as discussed in the previous section on thermal overstability. It is still hard to estimate theoretically the damping rate of a mode, particularly Γ_{conv} , because we need to treat “time-dependent” convection, for which theory has not yet been well established.

On the other hand, the dynamical damping rate consists of those due to the turbulent viscous dissipation, due to the leakage of waves from boundaries, and due to energy transfer to other modes by nonlinear interaction. That is,

$$\Gamma_{\text{dyn}} = \Gamma_{\text{visc}} + \Gamma_{\text{leak}} + \Gamma_{\text{coupl}}, \quad (54)$$

where the turbulent viscous damping and the damping due to the leakage of waves are supposed to be always positive but the term due to the mode coupling may be either positive or negative.

Since we assume in this picture the local balance between the stochastic driving by turbulent convection and the damping of the intrinsically stable p -modes, the energy of oscillation of the mode is then given by

$$E_q = G / \Gamma. \quad (55)$$

In the case of solar oscillation, the energy dissipation rate Γ is observationally determined by measuring the linewidth of the mode in the power spectrum, and the energy of oscillation of the mode E_q is also observationally determined by measuring the amplitude of the mode and by using theoretical eigenfunction ξ_q . In order to make progress in asteroseismology, however, we have to deduce rather theoretically the expected amplitude of oscillation by estimating G and Γ .

4.2 Estimate of acoustic power

The generation of acoustic noise by turbulence is known as the Lighthill mechanism (Lighthill 1952). Let us consider the acoustic power generated by a turbulent eddy of size λ_T having velocity v_λ , density ρ , and Mach number $M = v_\lambda/c$. According to the Lighthill theory, the total acoustic power generated by a turbulent eddy is given by

$$P \sim \frac{\rho v_\lambda^3}{\lambda_T} M^{2n+1}, \quad (56)$$

where n is the multipole index. Since the turbulence is thought to be subsonic, the lowest order multipole is the most effective. The lowest three multipoles are:

- monopole radiation ($n = 0$), which occurs when a parcel of gas executes periodical expansion and contraction,
- dipole radiation ($n = 1$), which occurs when an external force acts on the gas,
- quadrupole radiation ($n = 2$), which is emitted when turbulent Reynolds stresses act as the source term.

In the original Lighthill mechanism, the quadrupole radiation is the lowest order multipole generated by turbulence. In the following, we estimate the acoustic power based on the quadrupole radiation.

The characteristic time of the eddy (or the turn-over time) may be given by $\tau_\lambda \sim \lambda_T / v_\lambda$, and then the characteristic frequency of the eddy is given by $\sigma_{\text{eddy}} \sim v_\lambda / \lambda_T$. The characteristic frequency emitted by the quadrupole radiation is twice as high as this frequency, *i.e.*, $\sigma_Q \sim 2\sigma_{\text{eddy}}$ because the product of turbulent velocity components is involved in the Reynolds stress tensor $T_{ij} = \rho v_i v_j$ (Osaki 1990). Since the quadrupole radiation depends strongly on the Mach number, most of the acoustic radiation in the convection zone is generated near the top of the convection zone where the turbulent convective velocity attains its maximum value. We approximate this zone by a layer having a thickness of H at $r = r_s$ and the convective

velocity $v_{\max}$, and set $\lambda_T \sim H$ and $v_\lambda \sim v_{\max}$. Hence, $\sigma_Q \sim 2v_{\max}/H$, then the total acoustic power generated is given by

$$P(\sigma) \sim 4\pi R^2 H \epsilon(\sigma), \quad (57)$$

where R is the stellar radius and $\epsilon(\sigma)$ stands for the acoustic emissivity per unit volume and per unit frequency. If we assume the Kolmogoroff spectrum for the turbulent convective eddy, the acoustic emissivity is then given by (see, e.g., Goldreich and Kumar 1988)

$$\epsilon(\sigma) \sim \rho v_{\max}^2 M^5 \begin{cases} (\sigma/\sigma_Q)^2 & \text{for } \sigma \leq \sigma_Q \\ (\sigma/\sigma_Q)^{-7/2} & \text{for } \sigma \geq \sigma_Q. \end{cases} \quad (58)$$

The acoustic power per mode at the source is given by

$$G_{\text{source}} \sim P(\sigma) \frac{\Delta k}{k} \Delta \sigma, \quad (59)$$

where $\Delta k/k$ and $\Delta \sigma$ denote the horizontal wave number band-width for a particular non-radial mode and the frequency band-width, respectively. The horizontal wave-number band-width is estimated by assuming isotropic radiation as

$$\frac{\Delta k}{k} \sim \frac{1}{2(kR)^2} = \frac{1}{\sigma^2} \frac{c^2}{2R^2}. \quad (60)$$

Substituting equations (57), (58) and (60) into equation (59), we estimate the acoustic power *at the source*

$$G_{\text{source}} \sim \rho H^3 v_{\max}^2 M^3 \Delta \sigma \begin{cases} (\sigma/\sigma_Q)^0 & \text{for } \sigma \leq \sigma_Q \\ (\sigma/\sigma_Q)^{-5.5} & \text{for } \sigma \geq \sigma_Q. \end{cases} \quad (61)$$

It should be noted here that what we should estimate is the acoustic power *transmitted into the stellar acoustic cavity*. If the source layer is located within the cavity, the acoustic power in the cavity is G_{source} itself. On the other hand, if the source layer is located in the evanescent zone outside the acoustic cavity, the power is attenuated when it is transmitted into the cavity. Whether the source layer is inside the cavity or outside the cavity is dependent on the frequency under consideration. As shown in section 2.4, wave propagates as an acoustic wave only if $\sigma > \Phi_p(r)$ and $\sigma > \Phi_g(r)$. This condition is deduced to $\sigma > \sigma_{\text{ac}}(r)$ in the outer part of the star, since $\Phi_p(r) \simeq \sigma_{\text{ac}}(r)$ and $\Phi_g(r) \simeq \sigma_{\text{ac}}(r)$ there. Inversely, the wave is evanescent if $\sigma \leq \sigma_{\text{ac}}(r)$. In order to estimate the attenuation effect, let us assume that the $\sigma_{\text{ac}}(r)$ is a monotonically increasing function with r near the surface. Then, there is no attenuation effect for $\sigma > \sigma_{\text{ac}}(r_s)$, while the attenuation factor is $\exp[\int_{r_t}^{r_s} c^{-1}(1 - \sigma^2/\sigma_{\text{ac}}^2)^{1/2} dr]$ for $\sigma < \sigma_{\text{ac}}(r_s)$, where r_t is the radius at which $\sigma_{\text{ac}}(r_t) = \sigma$. Hence, in the case of $\sigma_{\text{ac}}(r_s) < \sigma_Q$, that is, if $v_{\max}/c(r_s) > 1/4$, the acoustic power *transmitted into the acoustic cavity*, G , is estimated as

$$G \sim \rho H^3 v_{\max}^2 M^3 \Delta \sigma \begin{cases} (\sigma/\sigma_Q)^{-5.5} & \text{for } \sigma \geq \sigma_Q \\ 1 & \text{for } \sigma_Q \geq \sigma \geq \sigma_{\text{ac}}(r_s) \\ \exp \left[- \int_{r_t}^{r_s} (1 - \sigma^2/\sigma_{\text{ac}}^2)^{1/2} \frac{dr}{c} \right] & \text{for } \sigma_{\text{ac}}(r_s) \geq \sigma. \end{cases} \quad (62)$$

On the other hand, if $\sigma_{\text{ac}}(r_s) > \sigma_Q$, that is, if $v_{\max}/c(r_s) < 1/4$, G is estimated as

$$G \sim \rho H^3 v_{\max}^2 M^3 \Delta \sigma \begin{cases} (\sigma/\sigma_Q)^{-5.5} & \text{for } \sigma \geq \sigma_{\text{ac}}(r_s) \\ \exp \left[- \int_{r_t}^{r_s} (1 - \sigma^2/\sigma_{\text{ac}}^2)^{1/2} \frac{dr}{c} \right] & \text{for } \sigma_{\text{ac}}(r_s) \geq \sigma. \end{cases} \quad (63)$$

It should be noted here that, for both the cases, for $\sigma > \sigma_{\text{ac}}(r_s)$, the wave leaks outward as a running wave since there is no reflection, and then the damping rate Γ_{leak} in equation (54) becomes large.

4.3 Numerical results

Expressions for the excitation rate have been derived analytically as discussed in the previous subsections. Evaluation of these expressions, however, depends on knowing the properties of convection, which are derived under appropriate approximations. Unfortunately, the theories of convection are still far from being complete. Both the traditional mixing length theory and the recent alternatively developed theories do not adequately describe the dynamics of convection and in addition contain free parameters and so lack predictive capability. In this situation, various attempts of hydrodynamical simulations have been made. Such simulations have succeeded in reproducing fairly well the properties of solar granulation, and then they are highly prospective for investigation of the stochastic excitation mechanism of p -mode oscillations. Using results of these numerical simulations of the near surface layer of a star it becomes possible to calculate the excitation rate of p -modes for the sun and other stars. Recent results are well summarized by Stein *et al.* (2004). The excitation rate of radial oscillations of stars near the main sequence from K to F and a subgiant K IV star has been calculated from numerical simulations of their surface convection zones. According to Stein *et al.* (2004), p -mode excitation increases with increasing effective temperature until envelope convection ceases in the F stars and also increases with decreasing gravity.

4.4 Comparison with observations

Detection of solar-like oscillations in late-type stars are vigorously attempted. In fact, detection of solar-like oscillations in those stars has been accomplished in the last few years, after decades of disappointing attempts (see Kurtz 2004). The bright star Procyon (α CMi) has been considered one of the best candidates for asteroseismology, on the basis of models and previous reports of p -modes detected in ground-based spectroscopy. A search for p -modes by nearly continuous photometric satellite-based observations was made by MOST (Matthews *et al.* 2004). However, the peak amplitudes are much smaller than the theoretical expectation and are inconsistent with the

ground-based observations of this star. More careful analyses are definitely needed to proceed with asteroseismology of solar-like stars.

Acknowledgements

I was honoured to have the occasion to attend the ARIES workshop which was held to celebrate the golden jubilee of the institution. I would like to express my sincere thanks to the SOC of the workshop for their kind invitation to the meeting and to all the members of ARIES for their warm hospitality during my stay in India.

References

- Baker, N., Kippenhahn, R. 1962, *Z. Astrophys.*, **54**, 114.
 Balmforth, N. J. 1992, *Monthly Notices Roy. Astron. Soc.*, **255**, 639.
 Balmforth, N. J., Cunha, M. S., Dolez, N., Gough, D. O., Vauclair, S. 2001, *Monthly Notices Roy. Astron. Soc.*, **323**, 362.
 Brickhill, A. J. 1991, *Monthly Notices Roy. Astron. Soc.*, **251**, 673.
 Chandrasekhar, S. 1964, *Astrophys. J.*, **139**, 664.
 Chaplin, W. J., Houdek, G., Elsworth, Y., Gough, D. O., Isaak, G. R., New, R. 2005, *Monthly Notices Roy. Astron. Soc.*, in press.
 Charpinet, S., Fontaine, G., Bassard, P. 2001, *Publ. Astron. Soc. Pacific*, **113**, 775.
 Cox, J. P. 1980, *Theory of Stellar Pulsation*, Princeton Univ. Press, Princeton.
 Cox, J. P., Cox, A. N., Olsen, K. H., King, D. S., Eilers, D. D. 1966, *Astrophys. J.*, **144**, 1038.
 Cowling, T. G. 1941, *Monthly Notices Roy. Astron. Soc.*, **101**, 367.
 Cowling, T. G. 1957, *Magnetohydrodynamics*, Interscience Tract, New York.
 Cunha, M. S. 2002, *Monthly Notices Roy. Astron. Soc.*, **333**, 47.
 Dziembowski, W. A., Moskalik, P., Pamyatnykh, A. A. 1993, *Monthly Notices Roy. Astron. Soc.*, **265**, 588.
 Eddington, A. S. 1926, *The Internal Constitution of the Stars*, Cambridge Univ. Press, Cambridge.
 Eisenfeld, J. 1969, *J. Math. Anal. Applic.*, **26**, 375.
 Fontaine, G., Wesemael, F. 1991, In *Proc. IAU Symp. No. 145, The photospheric abundance connection*, Reidel Publ. Co., Dordrecht, p. 421.
 Gautschi, A., Saio, H. 1993, *Monthly Notices Roy. Astron. Soc.*, **262**, 213.
 Gautschi, A., Saio, H. 1995, *Annual Rev. Astron. Astrophys.*, **33**, 75.
 Gautschi, A., Saio, H. 1996, *Annual Rev. Astron. Astrophys.*, **34**, 551.
 Goldreich, P., Keeley, D. A. 1977, *Astrophys. J.*, **212**, 243.
 Goldreich, P., Kumar, P. 1988, *Astrophys. J.*, **511**, 904.
 Goldreich, P., Wu, Y. 1999, *Astrophys. J.*, **326**, 462.
 Goldreich, P., Murray, N., Kumar, P. 1994, *Astrophys. J.*, **424**, 466.
 Gough, D. O., 1993, In: *Astrophysical Fluid Dynamics*, North-Holland, Amsterdam, p. 399.
 Houdek, G., Balmforth, N. J., Christensen-Dalsgaard, J., Gough, D. O. 1999, *Astron. Astrophys.*, **351**, 582.
 Iglesias, C. A., Rogers, F. J. 1993, *Astrophys. J.*, **412**, 752.
 Iglesias, C. A., Rogers, F. J. 1996, *Astrophys. J.*, **464**, 943.
 Kato, S. 1966, *Publ. Astron. Soc. Japan*, **18**, 374.
 Kiriakidis, M., El Eid, M. F., Glatzel, W. 1992, *Monthly Notices Roy. Astron. Soc.*, **255**, L1.
 Kurtz, D. W. 2004, *Solar Phys.*, **220**, 123.
 Ledoux, P. 1947, *Astrophys. J.*, **105**, 305.
 Ledoux, P., Walraven, Th. 1958, In *Handbuch der Physik, Bd. 51*, Springer-Verlag, Berlin, p. 353.
 Lighthill, M. J. 1952, *Proc. Roy. Soc. London A*, **211**, 564.
 Matthews, J. M., Kusching, R., Guenther, D. B., Walker, G. A. H., Moffat, A. F. J., Rucinski, S. M., Sasselov, D., Weiss, W. W. 2004, *Nature*, **340**, 51.
 Moskalik, P., Dziembowski, W. A. 1992, *Astron. Astrophys.*, **256**, L5.
 Osaki, Y. 1990, In: *Progress of Seismology of the Sun and Stars*, Springer-Verlag, Berlin, p. 75.

- Rogers, F. J., Iglesias, C. A. 1992, *Astrophys. J. Suppl.*, **79**, 507.
- Saio, H. 2005, *Monthly Notices Roy. Astron. Soc.*, in press.
- Samadi, R., Goupil, M.-J. 2001, *Astron. Astrophys.*, **370**, 136.
- Seaton, M. J., Yan, Y., Mihalas, D., Pradhan, A. K. 1994, *Monthly Notices Roy. Astron. Soc.*, **266**, 805.
- Shibahashi, H. 1983, *Astrophys. J.*, **275**, L5.
- Shibahashi, H., Osaki, Y. 1976, *Publ. Astron. Soc. Japan*, **28**, 199.
- Stein, R., Georgobiani, D., Trampedach, R., Ludwig, H.-G., Nordlund, A. 2004, *Solar Phys.*, **220**, 229.
- Unno, W., Osaki, Y., Ando, H., Saio, H., Shibahashi, H. 1989, *Nonradial Oscillations of Stars* (2nd edn), Univ. of Tokyo Press, Tokyo.
- Zhevakin, S. A. 1953, *Russian Astron. J.*, **30**, 161.

Helioseismology

H. M. Antia

*Tata Institute of Fundamental Research, Homi Bhabha Road, Mumbai 400 005, India.
e-mail: antia@tifr.res.in*

Abstract. The sun being the nearest star, seismic observations with high spatial resolution are possible, thus providing accurate measurement of frequencies of about half million modes of solar oscillations covering a wide range of degree. With these data helioseismology has enabled us to study the solar interior in sufficient detail to infer the large-scale structure and rotation of the solar interior. With the availability of high quality helioseismic data over a good fraction of a solar cycle it is also possible to study temporal variations in solar structure and dynamics. Some of these problems and recent results will be discussed.

Key words. Sun: oscillations—sun: interior—sun: rotation.

1. Introduction

The study of solar oscillations during the last three decades has provided a tool to study the solar interior in the same way as the study of seismic waves travelling through the earth have enabled the study of the earth's interior. The study of solar interior using oscillations has been referred to as helioseismology (Deubner & Gough 1984; Gough & Toomre 1991). Solar oscillations were discovered by Leighton *et al.* (1962) when they measured velocity at a point on the solar disk using the resulting Doppler shift. They found an oscillatory pattern with a period of around 5 minutes and hence these oscillations are also referred to as five-minute oscillations. The nature of these oscillations was not understood at that time and various theories were put forward to explain them. Ulrich (1970) and Leibacher & Stein (1971) suggested that these oscillations could be acoustic modes of oscillations of the entire sun which are trapped just below the photosphere. Subsequent observations with spatial resolution (Deubner 1975) confirmed this hypothesis as the power was found to be concentrated in a series of ridges in the k – ω diagram, where k is the spatial wavenumber and ω the temporal frequency of oscillations, in accordance with prediction from theoretical models. Once the nature of the modes was identified, it was realised that their properties are determined by the internal structure and dynamics of the sun. Hence, one can study the internal structure of the sun using these oscillations. It is now known that the observed oscillatory pattern on the solar surface arises from superposition of millions of independent modes of solar oscillations. Each of these modes provides an independent constraint on solar structure and dynamics.

Historically, stellar pulsation was first discovered on other stars where the amplitude of oscillations is much larger. Mira (*o Ceti*) is probably the first star which was established to be variable by Fabricius in 1596 and its oscillatory period was determined by

Holwarda in 1638, more than three hundred years before solar oscillations were discovered. While the luminosity in the visual band for this star varies by about an order of magnitude due to oscillations, for the sun the amplitude of oscillation of individual modes is of the order of the 10^{-6} of the solar luminosity. That is the main reason why solar oscillations were discovered only recently when appropriate technology to detect such low amplitude oscillations became available. Because of its proximity, the sun is the only star for which the disk can be easily resolved and hence it is possible to study a large number of modes, which may not be feasible for other stars.

The modes of solar oscillations can be characterized by 3 quantum numbers, the radial order n , the degree l and azimuthal order m . The radial order, n , is determined by the number of nodes along the radial direction, while l and m determine the horizontal structure of the eigenfunction which is defined by the spherical harmonic $Y_l^m(\theta, \phi)$ where θ is the colatitude and ϕ the longitude. Oscillations with $l \lesssim 3500$ have been detected on the sun. Early observations of solar oscillations had limited spatial and temporal resolution. It was soon recognised that in order to obtain high temporal resolution, one needs almost continuous observations covering long periods. This has been possible using a network of ground based observatories, e.g., the Global Oscillation Network Group (GONG) (Harvey *et al.* 1996) or from a suitably located satellite such as the Michelson Doppler Imager (MDI) instrument (Scherrer *et al.* 1995) on the Solar and Heliospheric Observatory (SOHO). The GONG network became operational in May 1995 and by now about 9 years of data have been analysed. The SOHO satellite was launched in December 1995 and the MDI instrument has been functioning since May 1996, except for two major breaks between July 1998 and February 1999, when the contact with the satellite was lost.

2. Seismic inferences of solar structure

If the sun were spherically symmetric, the frequencies $\nu_{n,l,m}$ of the modes of oscillations would be independent of m and these frequencies would be determined by the internal structure of the sun. The presence of rotation and magnetic field in the sun lifts the degeneracy of frequencies, but since the forces due to rotation and magnetic field are much smaller than those due to gravity and pressure gradient, we can treat rotation and magnetic field as small perturbations over the spherically symmetric structure. This gives rise to splitting of frequencies with the same n, l but different m . It is convenient to define the splitting coefficients by

$$\nu_{n,l,m} = \nu_{n,l} + \sum_j c_j^{n,l} P_j^{n,l}(m), \quad (1)$$

where, $\nu_{n,l}$ is the mean frequency of the multiplet, $c_j^{n,l}$ are the splitting coefficients and $P_j^{n,l}(m)$ are the set of orthogonal polynomials of degree j in m . The mean frequencies $\nu_{n,l}$ are determined by the horizontally averaged structure of the sun, while the splitting coefficients are determined by rotation, magnetic field and any other departure from spherical symmetry.

The mean frequencies $\nu_{n,l}$ of modes have been determined to very high accuracy. Figure 1 shows the l - ν diagram showing the frequencies as a function of l . The error bars in this figure show 1000σ errors in frequencies of each mode in the GONG data. For comparison, the frequencies of a standard solar model are also shown. It is clear that

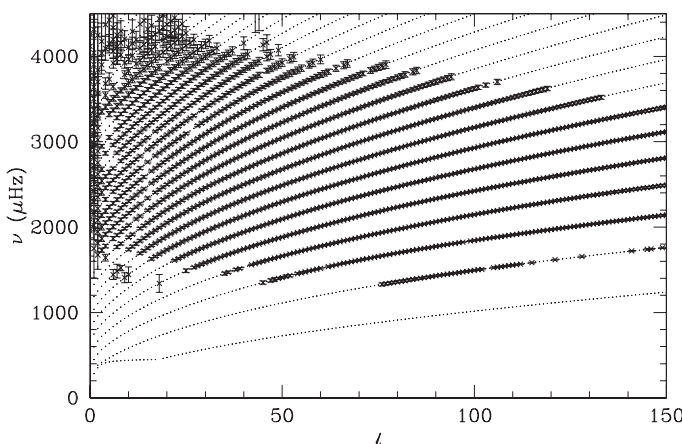


Figure 1. The l - ν diagram calculated using GONG data is compared with that for a standard solar model. The dots which have almost merged into a line represent frequencies of a solar model, while points with error bars are the observed frequencies. The error bars show 1000σ errors.

the model frequencies agree very well with the observed frequencies, thus giving us confidence in the solar model as well as the identification of observed modes with the acoustic modes of the sun. The various lines in this figure show the modes with different values of n starting with $n = 0$ as the lowest line. The $n = 0$ modes are referred to as the fundamental or f -mode and for large values of l these are essentially the surface gravity modes. The higher values of n correspond to the acoustic or p -modes where the pressure gradient provides the main restoring force. These are essentially sound waves propagating through the solar interior. Apart from these there is another series of modes known as gravity or g -modes with negative values of n . So far there is no reliable detection of these modes in the sun and hence these are not shown in Fig. 1.

Depending on the frequency and degree, the p -modes are trapped in different regions of the solar interior. As the sound waves travel inwards, the temperature and hence the sound speed increases and the waves are refracted away from the radial direction, until at some point they suffer a total internal reflection and are directed outwards. This defines the layer above which the mode is trapped. The critical layer depends on the angle of inclination to the radial direction at the surface, which is determined by the horizontal wavelength or the degree l . Thus modes with small values of l are closer to the vertical direction and hence penetrate to deeper layers before they get reflected outwards. While modes with large l are trapped in the layers immediately below the surface. The radial mode ($l = 0$) penetrates to the centre of the sun. This is an important property of acoustic modes which gives them the diagnostic power. Since different modes are trapped in different regions their frequencies are determined by the structure variation in the corresponding regions. Conversely, by studying the properties of a large number of acoustic modes we can infer the internal structure of the sun. A large number of inversion techniques have been developed for this purpose (Gough & Thompson 1991). Using these inversion techniques it is possible to determine the sound speed and density in the solar interior and the results are shown in Fig. 2. This figure shows the relative difference in sound speed and density between the sun and the standard solar model of Christensen-Dalsgaard *et al.* (1996).

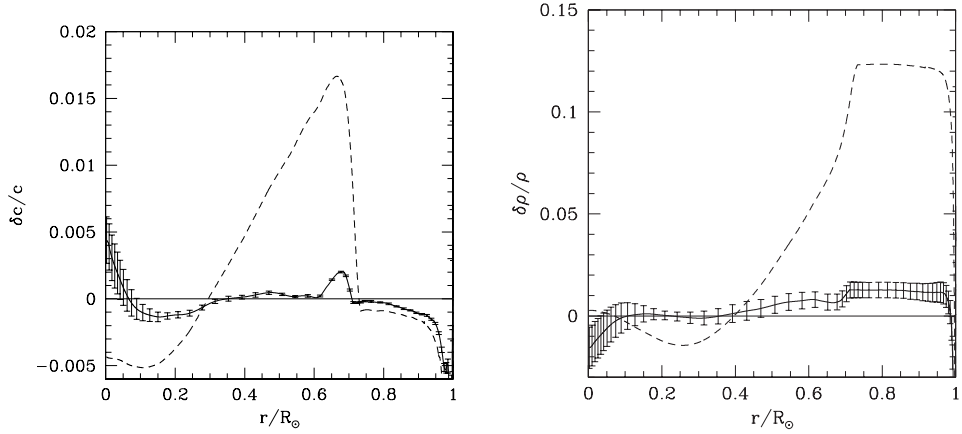


Figure 2. Relative difference in sound speed and density between the sun and the standard solar model of Christensen-Dalsgaard *et al.* (1996) is shown by *solid* lines with error bars. The *dashed* lines show the results obtained using a solar model with recent abundances.

From Fig. 2, it is clear that the sound speed and density in the standard solar model is close to that in the sun. This along with other considerations led to the conclusion that the low solar neutrino fluxes observed in the early experiments are due to neutrino oscillations (Bahcall *et al.* 2001 and references therein). Recently, these neutrino oscillations have been confirmed by neutrino detectors at the Sudbury Neutrino Observatory (Ahmad *et al.* 2002). This was regarded as a significant achievement of solar models and helioseismology. However, soon after this, the agreement between the solar model and the seismically inferred structure deteriorated when Asplund *et al.* (2004a, b) found that the oxygen abundance in the solar photosphere should be reduced by a factor of 1.5. Along with oxygen the abundances of other elements were also reduced. This reduces Z/X in the solar envelope from 0.023 (Grevesse & Sauval 1998) to 0.0165 (Asplund *et al.* 2004b), thus reducing the opacity of the solar material. This has introduced significant differences between the standard solar model and the seismically inverted structure profiles (Basu & Antia 2004; Bahcall *et al.* 2005) as well as in the depth of the convection zone (Bahcall & Pinsonneault 2004). Figure 2 also shows the difference in sound speed and density between a solar model constructed with revised abundances and the sun. It is clear that the differences are almost an order of magnitude larger than those with the earlier model using older abundances. The model with revised abundances also has a shallow convection zone. In order to get the correct depth of the convection zone, it is necessary to increase the opacity by about 20% as compared to the standard OPAL (Iglesias & Rogers 1996) values near the base of the convection zone (Basu & Antia 2004; Bahcall *et al.* 2004; Guzik & Watson 2005). It is unlikely that errors in theoretically computed opacities are so large. Recent independent computations by the OP project (Badnell *et al.* 2004) find a difference of only 2% between OP and OPAL opacities near the base of the convection zone. Increased diffusion of helium and heavy elements below the convection zone also does not solve the problem as in that case the resulting helium abundance in the solar envelope is much less than the seismically inferred value (Basu & Antia 2004).

The fact that a solar model using older abundances is remarkably close to the seismically inferred structure suggests that the revised abundances need to be checked

independently. The revised abundances have been obtained using improved 3D hydrodynamic model atmosphere. Nevertheless, there could be some systematic errors due to inadequate treatment of turbulence or other effects. Antia & Basu (2005) have found that among all heavy elements the abundances of O, Ne and Fe are most important in determining the opacity near the base of the convection zone. Of these elements, the abundance of Ne has not been determined from photospheric lines and hence there could be larger uncertainties in its determination. If the abundance of Ne is increased by a factor of 2.5 as compared to the recently reduced value, it can compensate for much of the reduction in oxygen abundance (Antia & Basu 2005). It is not clear if this much increase in Ne abundance is permissible. It is clear that more work is required to verify the abundance determination as well as opacity calculations to resolve this discrepancy.

3. Rotation rate in the solar interior

From the observed rotational splitting of modes it is possible to determine the rotation rate in the solar interior (Thompson *et al.* 1996; Schou *et al.* 1998). Inversions of observed splitting coefficients have shown that the observed differential rotation at the solar surface continues throughout the convection zone, but near the base of the convection zone there is a transition to rotation rate that is independent of latitude. In the radiative interior the rotation rate is roughly constant. The transition region where the rotation rate changes from differential rotation to solid body like rotation is referred to as the tachocline (Spiegel & Zahn 1992). This is a region with strong shear and is generally believed to be the layer where solar dynamo is operating. This shear layer is also believed to cause some mixing in the tachocline region (Richard *et al.* 1996; Brun *et al.* 1999).

With the accumulation of seismic data over the last nine years it is possible to study the temporal variation in the solar structure and rotation rate. This may help us in understanding the operation of solar dynamo. The frequencies of solar oscillations are known to vary with solar activity and in fact, the variation is well correlated with solar activity (Libbrecht & Woodard 1990; Howe *et al.* 1999; Bhatnagar *et al.* 1999). These frequency differences can be inverted to study the variation in solar structure with solar cycle. These inversions show that almost all of the observed variation in solar oscillation frequencies is due to variation in surface layers and there is no significant variation in the solar interior. Basu and Mandel (2004) have found a marginally significant variation in the He II ionisation region close to the solar surface. However, inversion of rotational splittings show a significant temporal variation in the rotation rate in the solar interior (Howe *et al.* 2000; Antia & Basu 2000) which is similar to the observed torsional oscillations (Howard & LaBonte 1980) at the solar surface.

To study the temporal variations in the rotation rate we can define the residual rotation rate:

$$\delta\Omega(r, \theta, t) = \Omega(r, \theta, t) - \langle \Omega(r, \theta, t) \rangle, \quad (2)$$

where the angular brackets denote averaging over time. It is often more convenient to show the zonal flow velocity $\delta v_\phi = \delta\Omega r \cos \theta$ where θ is the latitude. Figure 3 shows the contours of constant zonal flow velocity at $r = 0.98R_\odot$ and at $\theta = 15^\circ$. These figures show a clear pattern with bands of faster (or slower) than average velocity

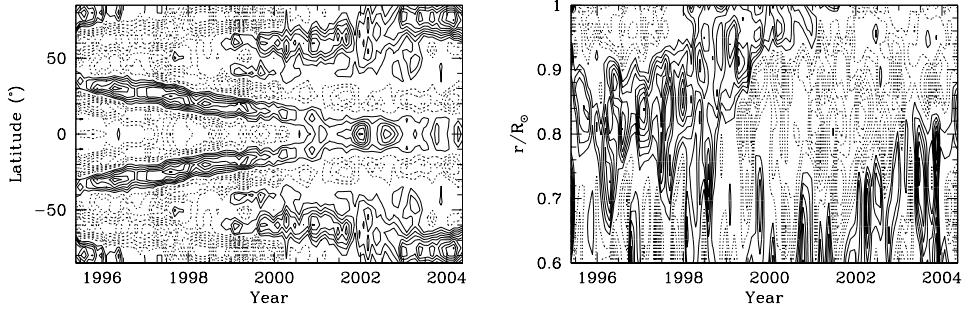


Figure 3. Contours of constant zonal flow velocity are shown as a function of time and latitude at $r = 0.98R_{\odot}$ (left panel) and as a function of time and radius at a latitude of 15° (right panel). The continuous lines show positive values and the dotted lines show negative values. The contour interval is 1 m s^{-1} .

moving with time. Further, these bands move towards the equator at low latitudes (Howe *et al.* 2000; Antia & Basu 2000) while at high latitudes these bands move towards the poles (Antia & Basu 2001). From the figure for $\theta = 15^{\circ}$ it can be seen that the bands are rising upwards with time and they continue almost to the base of the convection zone. Thus the zonal flow pattern continues throughout the convection zone.

4. The solar radius

The standard value of solar radius, $R_{\odot} = (695.99 \pm 0.07) \text{ Mm}$ was first obtained by Auwers (1891). Since then a number of improved measurements have been made giving different results. Because the sun is not a solid object there is some ambiguity in the definition of solar surface and hence its radius. Conventional measurements use the point of inflexion in the intensity profile as the location of surface, while the solar models use a level where the optical depth (measured from outside) is of an order of unity. Using detailed atmospheric models Brown & Christensen-Dalsgaard (1998) have shown that there would be a difference of about 500 km between the two definitions of the solar radius, with the model definition being smaller. The solar oscillation frequencies are very sensitive to the solar radius and hence can be used to determine the solar radius. In particular, the f -modes, which are essentially surface gravity modes are well suited for this purpose. The frequencies of f -modes are essentially independent of stratification and are approximately given by the dispersion relation $\omega^2 \approx gk = GM\sqrt{l(l+1)}/r^3$ where g is the acceleration due to gravity at the surface, $k = \sqrt{l(l+1)}/r$ is the horizontal wave number, G is the gravitational constant. By comparing the observed frequencies with that in a solar model with given radius it is possible to estimate the radius (Schou *et al.* 1997; Antia 1998). Using observed frequencies from GONG and MDI data it is found that the solar model radius should be reduced by 200–300 km from the standard value. The exact value depends on the modelling of near surface layers in the solar model (Tripathy & Antia 1999). These systematic errors arising from the model are expected to be independent of time and hence the f -mode frequencies may be used to detect temporal variations in the solar radius at a level of 1 km.

There are conflicting reports about temporal variation in solar radius, with claims ranging from 0 to 700 km (Laclare *et al.* 1996 and references therein). Recent observations from MDI have put an upper limit of 5 km on possible temporal variation in solar radius during the current solar cycle (Kuhn *et al.* 2004). It can be easily seen that a variation in solar radius by δR will change the gravitational potential energy by $(GM_{\odot}^2/R_{\odot}^2)\delta R \approx 5 \times 10^{42}\delta R$ ergs, if δR is in km. If this variation is over a solar cycle of 11 years (3.5×10^8 sec) the energy is released or absorbed at the rate of $1.4 \times 10^{34}\delta R$ ergs per second. Thus even if $\delta R = 1$ km the rate of energy variation is more than the solar luminosity. Hence it is clear that if there is any significant variation in solar radius, it must be confined to outer layers which have little mass. Even if the radius variation extends till the base of the convection zone, the energy will be released or absorbed at a rate that is 2% of the above estimate, which is still quite large.

Thus it would be interesting to study these temporal variations using f -mode frequencies. Unfortunately, these studies have also given conflicting results with variations between 0 and 5 km (Dziembowski *et al.* 1998, 2000, 2001; Antia *et al.* 2000, 2001; Antia 2003). The dispersion in the results is because the actual variation in f -mode frequencies is more complicated with at least two independent components (Antia *et al.* 2001). One of the time varying components is found to have a period of almost exactly 1 year and is most likely to be an artifact in data due to the orbital motion of the earth. The second component appears to be correlated with solar activity but it has a steep dependence on degree l and hence is unlikely to be due to radius variation. Dziembowski *et al.* (2001, henceforth DGS) tried to decompose the f -mode frequency variation into two components, one arising from radius variation and another from some variation in the surface layer, which is expected to scale inversely with the mode inertia. Thus they assume

$$\Delta v_{l,0} = -\frac{3}{2} \frac{\Delta R}{R} v_{l,0} + \frac{\Delta \gamma}{I_{l,0}}, \quad (3)$$

where $\Delta v_{l,0}$ is the variation in f -mode ($n = 0$) frequency, ΔR is the variation in radius, $I_{l,0}$ is the mode inertia and $\Delta \gamma$ measures the variation in the surface contribution. This expression does not account for the oscillatory component with a period of one year. Nevertheless, DGS computed ΔR and $\Delta \gamma$ for each set of observed frequencies from MDI instrument and concluded that the radius is decreasing at the rate of (1.5 ± 0.3) km per year during the rising phase of solar activity. Figure 4 shows the result of the same exercise using an extended data set that is now available. It can be seen the χ^2 per degree of freedom in these fits are fairly large and it is clear that the data cannot be fitted by a simple expression given by equation (3). Further, the resulting variation in $\Delta \gamma$ is correlated to the solar activity, while radius does appear to decrease with time. Both these quantities show variation with a time period of 1 year, which has not been removed from the data. Apart from this oscillatory variation it appears that the radius has suddenly decreased by about 4 km around 1999, rather than a gradual decrease as claimed by DGS. If this variation is real then it would be difficult to explain. This time period coincides with the period when contact with SOHO satellite was lost and it is quite likely to be due to some systematic errors introduced during the recovery of the satellite.

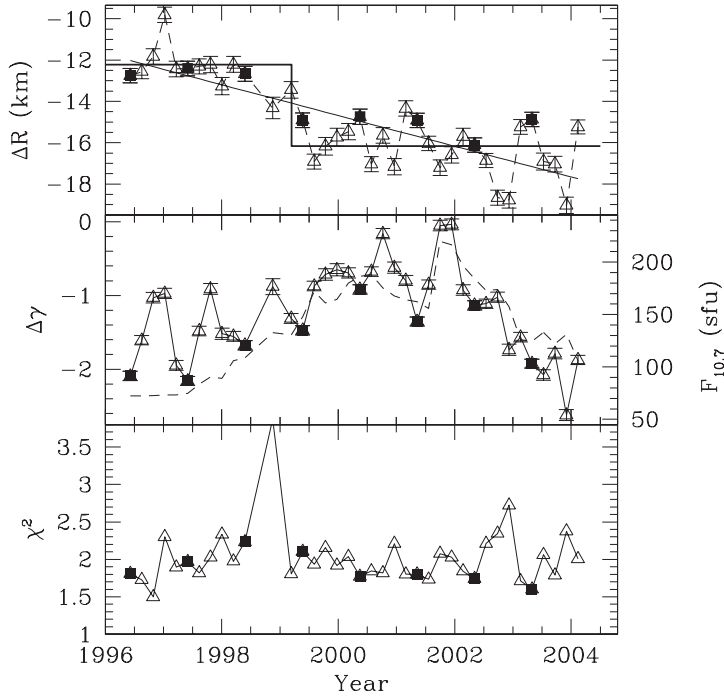


Figure 4. The estimated variation in the solar radius, ΔR , and the surface term, $\Delta\gamma$ from f -mode frequencies, obtained by fitting equation 3 to frequency difference between a given MDI set and a solar model. The χ^2 per degree of freedom for each set is shown in the lowest panel. In each panel the filled squares are the results for data sets at an interval of 360 days for which the fit is relatively good. The solid line in the top panel is a straight line fit to all points, similar to that obtained by DGS. The heavy line shows a step function fit to all points with discontinuity at 1999.2. The dashed line in the middle panel shows the 10.7 cm radio flux on a scale shown at the right.

References

- Ahmad, Q. R. *et al.* 2002, *Phys. Rev. Lett.*, **89**, 011301.
 Antia, H. M. 1998, *Astron. Astrophys.*, **330**, 336.
 Antia, H. M. 2003, *Astrophys. J.*, **590**, 567.
 Antia, H. M., Basu, S. 2000, *Astrophys. J.*, **541**, 442.
 Antia, H. M., Basu, S. 2001, *Astrophys. J.*, **559**, L67.
 Antia, H. M., Basu, S. 2005, *Astrophys. J.*, **620**, L129.
 Antia, H. M., Basu, S., Pintar, J., Pohl, B. 2000, *Solar Phys.*, **192**, 459.
 Antia, H. M., Basu, S., Pintar, J., Schou, J. 2001, In *Helio- and Asteroseismology at the Dawn of the Millennium*, ESA SP-464, p. 27.
 Asplund, M., Grevesse, N., Sauval, A. J., Allende Prieto, C., Kiselman, D. 2004a, *Astron. Astrophys.*, **417**, 751.
 Asplund, M., Grevesse, N., Sauval, A. J. 2004b, In *Cosmic abundances as records of stellar evolution and nucleosynthesis*, ASP Conf. Series, (in press) (astro-ph/0410214).
 Auwers, A. 1891, *Astron. Nach.*, **128**, 361.
 Badnell, N. R., Bautista, M. A., Butler, K., Delahaye, F., Mendoza, C., Palmeri, P., Zeippen, C. J., Seaton, M. J. 2004, astro-ph/0410744.
 Bahcall, J. N., Pinsonneault, M. H. 2004, *Phys. Rev. Lett.*, **92**, 121301.
 Bahcall, J. N., Pinsonneault, M. H., Basu, S. 2001, *Astrophys. J.*, **555**, 990.
 Bahcall, J. N., Serenelli, A. M., Pinsonneault, M. H. 2004, *Astrophys. J.*, **614**, 464.
 Bahcall, J. N., Basu, S., Pinsonneault, M. H., Serenelli, A. M. 2005, *Astrophys. J.*, **618**, 1049.

- Basu, S., Antia, H. M. 2004, *Astrophys. J.*, **606**, L85.
- Basu, S., Mandel, A. 2004, *Astrophys. J.*, **617**, L155.
- Bhatnagar, A., Jain, K., Tripathy, S. C. 1999, *Astrophys. J.*, **521**, 885.
- Brown, T. M., Christensen-Dalsgaard, J. 1998, *Astrophys. J.*, **500**, L195.
- Brun, A. S., Turck-Chièze, S., Zahn, J.-P. 1999, *Astrophys. J.*, **525**, 1032.
- Christensen-Dalsgaard, J. *et al.* 1996, *Science*, **272**, 1286.
- Deubner, F.-L. 1975, *Astron. Astrophys.*, **44**, 371.
- Deubner, F.-L., Gough, D. O. 1984, *Ann. Rev. Astron. Astrophys.*, **22**, 593.
- Dziembowski, W. A., Goode, P. R., DiMauro, M. P., Kosovichev, A. G., Schou, J. 1998, *Astrophys. J.*, **509**, 456.
- Dziembowski, W. A., Goode, P. R., Kosovichev, A. G., Schou, J. 2000, *Astrophys. J.*, **537**, 1026.
- Dziembowski, W. A., Goode, P. R., Schou, J. 2001, *Astrophys. J.*, **553**, 897 (DGS).
- Gough, D. O., Thompson, M. J. 1991, In *Solar interior and atmosphere*, Space Science Series, University of Arizona Press, p. 519.
- Gough, D. O., Toomre, J. 1991, *Ann. Rev. Astron. Astrophys.*, **29**, 627.
- Grevesse, N., Sauval, A. J. 1998, In *Solar composition and its evolution – from core to corona*, Kluwer, Dordrecht, p. 161.
- Guzik, J. A., Watson, L. S. 2005, In *Helio- and Asteroseismology: Towards a golden future*, ESA SP-559, (astro-ph/0501530), p. 456.
- Harvey, J. W. *et al.* 1996, *Science*, **272**, 1284.
- Howard, R., LaBonte, B. J. 1980, *Astrophys. J.*, **239**, L33.
- Howe, R., Komm, R., Hill, F. 1999, *Astrophys. J.*, **524**, 1084.
- Howe, R., Christensen-Dalsgaard, J., Hill, F., Komm, R. W., Larsen, R. M., Schou, J., Thompson, M. J., Toomre, J. 2000, *Astrophys. J.*, **533**, L163.
- Iglesias, C. A., Rogers, F. J. 1996, *Astrophys. J.*, **464**, 943.
- Kuhn, J. R., Bush, R. I., Emilio, M., Scherrer, P. H. 2004, *Astrophys. J.*, **613**, 1241.
- Laclare, F., Delmas, C., Coin, J. P., Irbah, A. 1996, *Solar Phys.*, **166**, 211.
- Leibacher, J., Stein, R. F. 1971, *Astrophys. Lett.*, **7**, 191.
- Leighton, R. B., Noyes, R. W., Simon, G. W. 1962, *Astrophys. J.*, **135**, 474.
- Libbrecht, K. G., Woodard, M. F. 1990, *Nature*, **345**, 779.
- Richard, O., Vauclair, S., Charbonnel, C., Dziembowski, W. A. 1996, *Astron. Astrophys.*, **312**, 1000.
- Scherrer, P. H. *et al.* 1995, *Solar Phys.*, **162**, 129.
- Schou, J., Kosovichev, A. G., Goode, P. R., Dziembowski, W. A. 1997, *Astrophys. J.*, **489**, L197.
- Schou, J. *et al.* 1998, *Astrophys. J.*, **505**, 390.
- Spiegel, E. A., Zahn, J.-P. 1992, *Astron. Astrophys.*, **265**, 106.
- Thompson, M. J. *et al.* 1996, *Science*, **272**, 1300.
- Tripathy, S. C., Antia, H. M. 1999, *Solar Phys.*, **186**, 1.
- Ulrich, R. K. 1970, *Astrophys. J.*, **162**, 993.

Excitation of Solar-like Oscillations: From PMS to MS Stellar Models

R. Samadi¹, M.-J. Goupil¹, E. Alecian¹, F. Baudin², D. Georgobiani³,
R. Trampedach⁴, R. Stein⁵ & Å. Nordlund⁶

¹*Observatoire de Paris, 5 place J. Janssen, 92195 Meudon, France.*

²*Institut d'Astrophysique Spatiale, Orsay, France.*

³*Center for Turbulence Research, Stanford University NASA Ames Research Center, Moffett Field, USA.*

⁴*Research School of Astronomy and Astrophysics, Mt. Stromlo Observatory, Weston, Australia.*

⁵*Michigan State University, East Lansing, USA.*

⁶*Niels Bohr Institut, Copenhagen, Denmark.*

Abstract. The amplitude of solar-like oscillations results from a balance between excitation and damping. As in the sun, the excitation is attributed to turbulent motions that stochastically excite the p modes in the uppermost part of the convective zone. We present here a model for the excitation mechanism. Comparisons between modeled amplitudes and helio and stellar seismic constraints are presented and the discrepancies discussed. Finally the possibility and the interest of detecting such stochastically excited modes in pre-main sequence stars are also discussed.

Key words. Turbulence—convection—oscillations—excitation—sun, stars: α Cen A—stars: main and pre-main sequence stars.

1. Introduction

In the past approximately five years, solar-like oscillations have been detected in several intermediate massive stars lying on the main sequence (e.g., Procyon, α Cen A and B, β Hydri, β Vir, HD 49933, etc.), in early post-main sequence stars (e.g., η Boo) and also in red giant stars (ξ Hya, ϵ Oph, η Ser). We refer the reader to the recent review by Kjeldsen & Bedding (2004).

Those oscillations are stochastically excited by turbulent motions in the uppermost part of the convective zone. This mechanism has been modeled by several authors in different ways (e.g., Goldreich & Keeley 1977; Christensen-Dalsgaard & Frandsen 1983; Balmforth 1992; Goldreich *et al.* 1994; Samadi & Goupil 2001). Those approaches differ from each other by the nature of the assumed excitation sources, the assumed simplifications and approximations and also by the way turbulent convection is described (see the review by Stein *et al.* 2004).

Among the different theoretical approaches, that by Samadi & Goupil (2001, Paper I hereafter) includes a detailed treatment of the properties of turbulent convection. This formalism then offers the advantage of testing different assumptions about turbulent convection in stars. We briefly present this formalism in section 2 (see also the detailed summary by Samadi 2001). Section 3 explains how the mode excitation rates may be

computed – according to this formalism – from 1D stellar models or 3D simulation of stellar convection.

Provided that measurements of the oscillation amplitudes and life-times (or line-widths) are available it is possible to compute the rate P at which energy is injected into the modes and hence to derive constraints on the models of stochastic excitation. The way those constraints are derived from seismic observations is summarized in section 4.

Such measurements have been available for the sun for several years (e.g., Chaplin *et al.* 1997, 1998; Libbrecht 1988). Based on the approach of Paper I and constraints from a 3D simulation of the sun, Samadi *et al.* (2003b, hereafter Paper IV) have found a rather good agreement between predicted excitation rates P and excitation rates inferred from the helioseismic observations by Chaplin *et al.* (1998), except at high frequency where the discrepancy is substantial. However, recently, Baudin *et al.* (2005) have investigated in detail, the difficult problem of inferring solar p mode excitation rates from the helioseismic data. Large differences between the new constraints obtained by Baudin *et al.* (2005) and those derived from Chaplin *et al.*'s (1998) are found. The computed excitation rates P are found to be in even better agreement with the new constraints.

Recently Bedding *et al.* (2004) have derived – from recent seismic observations by Butler *et al.* (2004) – oscillation amplitudes and also averaged estimates of the oscillation life-times for α Cen A. Those data enable us to constrain P for that star. The approach adopted in the case of the sun has thus been extended to the case of α Cen A and large discrepancies have been found for this star between predicted P and “observed” P (see Samadi *et al.* 2004).

For other stars, we may not yet derive observational constraints on P but only on the mode amplitudes (mode linewidths are then taken from theoretical calculations). Theoretical calculations performed in the past approximately five years, resulted in an *overestimation* of the amplitudes of the solar-like oscillations detected in stars hotter and more massive than the sun such as η Bootis, Procyon, ξ Hydrae (see Kjelson & Bedding 2001; Houdek & Gough 2002). This overestimation of the mode amplitudes may be attributed either to an overestimation of the excitation rates or an underestimation of the damping rates. In turn, any overestimation of the excitation rates may be attributed either to the excitation model *itself* or to the underlying convection model.

Numerical 3D simulations enable one to compute directly P for stars with various temperatures and luminosities (see Nordlund & Stein 2001). Hence, using this method, Samadi *et al.* (2005a, Paper V) have computed, for seven 3D simulations of outer layers of main sequence stars – for which solar-like oscillations are expected – the rates at which energy is expected to be injected into p modes. Comparison between these 3D calculations and those obtained using the theoretical model of stochastic excitation then provide a test of the validity of that theoretical model across the HR diagram. This test performed by Paper V is summarized in section 6.1.

Solar-like oscillations have so far been detected in main sequence stars as well as in red giant stars. However Samadi *et al.* (2002) predicted that one may also expect solar-like oscillations with large amplitudes in δ Scuti stars. As mentioned in that paper, there would be advantages in observing simultaneously opacity-driven and stochastically excited modes in the same δ Scuti star.

Pre-Main Sequence stars (PMS hereafter) are young objects on their way to the main sequence (MS hereafter). Prior to their arrival to the Zero Age Main Sequence

(ZAMS), their outer layers have the same properties as their MS counterpart in the HR diagram. For this reason, we expect stochastically excited modes in such objects. Detecting such oscillations would be a valuable way of probing the modeling of such active and very fast rotating young objects.

We have computed, for a representative set of Herbig Ae stars and T-Tauri stars, the rates at which energy is expected to be injected into the stable p modes of such stars. The results of this new investigation are presented in section 6.2.

Computation of the amplitudes of solar-like oscillations in terms of Doppler velocity (V) requires – in addition to P – the calculation of the mode damping rates η . Such calculations must be performed with a non-adiabatic pulsation code that takes into account not only the modulation of the radiative flux by the pulsation but also the interaction between pulsation and convection. Modeling this interaction is difficult. Several attempts have been made such as Gough’s (1977) time-dependent model of convection, Gabriel’s formalism (see e.g., Grigahcène *et al.* 2005) and Xiong’s approach (see e.g., Xiong *et al.* 1998).

Using Gough’s approach, Samadi *et al.* (2001b) have calculated η – as in Houdek *et al.* (1999) – for a set of solar-like oscillating stars lying on the main sequence. Finally from that calculation and predicted values of P , we have computed $V_{\max}$, the maximum of V , across the main sequence for different assumptions (Paper V). It was found that, provided that the improved model of Paper IV is adopted, computed $V_{\max}$ are closer to the observed one than previous calculations. Section 7.1 presents this result.

Finally we extrapolate, in section 7.2, the calculations of $V_{\max}$ – obtained for main sequence stars – for pre-main sequence stars and we conclude whether or not solar-like oscillations may be detected in such stars with current or future instruments devoted to stellar seismology.

Section 8 is devoted to a discussion of the current status of modeling the excitation of solar-like oscillations.

2. The theoretical model of stochastic excitation

The theoretical model of stochastic excitation considered here is basically that of Paper I. In this model, two sources are associated with the driving of the resonant modes of the stellar cavity: the first one is related to the Reynolds stress tensor and as such represents a mechanical source of excitation. The second one is caused by the advection of the turbulent fluctuations of entropy by the turbulent motions (the so-called “entropy source term”) and as such represents a thermal source of excitation.

The problem of the mode driving is solved in the framework of Goldreich & Keeley (1977)’s approach: the method consists of solving first the propagation of the *adiabatic* modes in the stellar cavity *without* the presence of turbulent convection. The solutions of this problem are the well-known adiabatic (real) displacement eigenvectors, that is: $\vec{\xi}(\vec{r}, t) = \vec{\xi}(\vec{r})e^{-i\omega_0 t}$ where ω_0 is the oscillation eigenfrequency. We then add the turbulence and search the general solution for the mode velocity, $\vec{v}_{\text{osc}}$, in the form:

$$\vec{v}_{\text{osc}} = \frac{1}{2}(i\omega_0 A(t) \vec{\xi}(\vec{r}) e^{-i\omega_0 t} + cc), \quad (1)$$

where cc means complex conjugate and $A(t)$ is an instantaneous amplitude that accounts for the driving by turbulent convection.

From the perturbed equation of continuity and motion and with the help of several simplifications, one derives a general expression for the mean square of A (see details in Paper I):

$$\langle |A|^2 \rangle = \frac{1}{8\eta(\omega_0 I)^2} (C_R^2 + C_S^2), \quad (2)$$

where η is the mode damping rate, $I \equiv \int_V \rho_0 d^3x (\vec{\xi}^* \cdot \vec{\xi})$ is the mode inertia, C_R^2 and C_S^2 are the turbulent Reynolds stress and entropy contributions respectively:

$$C_R^2 = \frac{64}{15} \pi^3 \int_0^M dm \rho_0 \left(\frac{d\xi_r}{dr} \right)^2 \int_0^\infty dk \int_{-\infty}^\infty d\omega \frac{E^2(k)}{k^2} \chi_k(\omega_0 + \omega, r) \chi_k(\omega, r), \quad (3)$$

$$C_S^2 = \frac{16}{3} \frac{\pi^3}{\omega_0^2} \int d^3x_0 (\alpha_s g_r)^2 g_r \int dk \int_{-\infty}^{+\infty} d\omega \frac{E_s(k) E(k)}{k^2} \chi_k(\omega_0 + \omega) \chi_k(\omega). \quad (4)$$

In equation (3), $E(k)$ is the turbulent kinetic energy spectrum, $\chi_k(\omega)$ is the eddy time-correlation function in the frequency space, ρ_0 is the mean density. In equation (4), $E_s(k)$ is the turbulent spectrum of the entropy fluctuations, $g_r(\xi_r, r)$ is a function that involves the first and the second derivatives of ξ (see Samadi *et al.* (2003a, Paper III hereafter)) and finally $\alpha_s \equiv (\partial p / \partial s)_\rho$ where p is the pressure and s the entropy.

The time-averaged mode energy is given by

$$E = \int_0^M dm \langle v_{\text{osc}}^2 \rangle = \frac{1}{2} \langle |A|^2 \rangle I \omega_0^2, \quad (5)$$

where $\langle |A|^2 \rangle$ is the mean square amplitude as given by equation (2). The energy supply per unit time into each mode is given by (see Samadi *et al.* 2001, hereafter Paper II):

$$P \equiv \frac{d}{dt} E = 2\eta E = \eta \langle |A|^2 \rangle I \omega_0^2. \quad (6)$$

The kinetic energy spectrum $E(k)$ is normalized as

$$\frac{1}{2} \langle u^2(\vec{x}_0) \rangle \equiv \int_0^\infty dk E(k) = \frac{1}{2} \Phi w^2, \quad (7)$$

where w is the rms of the vertical component of the velocity and Φ is a mean anisotropy factor introduced by Gough (1977) to take into account anisotropy effects at large scales (see equation (2) of Paper IV). The turbulent pressure is defined as usual as $p_t \equiv \rho_0 w^2$. Accordingly the driving by the Reynolds term is the largest where p_t is the largest, that is generally at the top of the convective envelope.

As for $E(k)$, the entropy spectrum $E_s(k)$ is normalized as:

$$\frac{1}{2} \langle s_t^2(\vec{x}_0) \rangle = \int_0^\infty dk E_s(k) = \frac{1}{2} \tilde{s}^2, \quad (8)$$

where $\tilde{s}$ is the rms of the entropy fluctuations. One shows that $\tilde{s}^2 = F_c^2 / (\rho_0 T_0^2 p_t)$ where F_c is the convective flux and T_0 the mean temperature (see Samadi *et al.* 2005b). Accordingly the driving by the so-called entropy source term is the largest where $F_c / (\rho_0 T_0)$ is the largest, that is at the top of the convective envelope where the density and temperature drop rapidly.

3. Methods for the calculation of the p mode excitation rates

The rate P at which energy is injected per unit time into a mode is calculated according to the set of equations (6), (2), (3) and (4). The calculation thus requires the knowledge of quantities that are of three different types:

- (1) Quantities that are related to the oscillation modes: the eigenfunctions (ξ_r) and associated eigenfrequencies (ω_0).
- (2) Quantities that are related to the spatial and time averaged properties of the medium: ρ_0 , α_s , w , $\tilde{s}$ and Φ .
- (3) Quantities that contain information about spatial and temporal correlations of the convective fluctuations: $E(k)$, $E_s(k)$ and $\chi_k(\omega)$.

Eigenfrequencies and eigenfunctions (in (1) above) are computed with an adiabatic pulsation model using a 1D stellar model. The spatial and time averaged quantities in (2) may be obtained either from a 3D simulation of the stellar surface or from a 1D stellar model. In the latter case, the calculation of w and $\tilde{s}$ relies on the mixing-length formulation of convection and introduces free parameters (such as the mixing-length parameter and Φ). Finally, for the quantities in (3), one may either obtain constraints from a 3D simulation or assume some models such as those investigated in Paper II. In the latter case, new adjustable parameters must generally be introduced as the scale length of the most energetic eddies and the parameter λ , which controls our lack of precise knowledge of the eddy time-correlations (see Paper IV).

The results presented here are based on different strategies and assumptions for the calculation of the quantities in (1), (2) and (3).

4. Inferring the mode excitation rates from seismic observations

The mode excitation rates may be inferred from the observations according to the relation:

$$P_{\text{obs}}(\omega_0) = 2\pi \Gamma_v \mathcal{M} v_s^2(\omega_0), \quad (9)$$

where $\mathcal{M} \equiv I/\xi_r^2(h)$ the mode mass, h the height above the photosphere where oscillations are measured, $\Gamma_v = \eta/\pi$ the mode linewidth at half maximum (in Hz) and v_s the rms of the mode surface velocity.

As pointed out by Baudin *et al.* (2005), the layer (h) where the mode mass is evaluated must be properly estimated to derive correct values of the excitation rates. Indeed solar seismic observations in Doppler velocity are usually measured from a given spectral line. The layer where oscillations are measured then depends on the height where the line is formed. Different instruments use different solar lines and

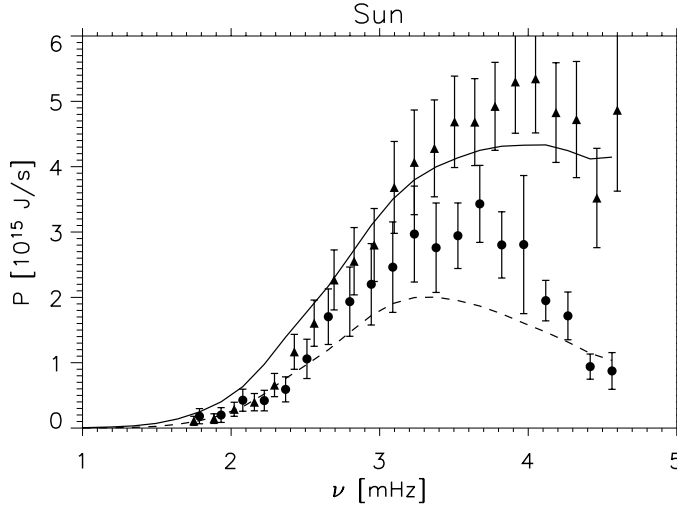


Figure 1. Rates P at which energy is injected into the solar modes. Filled circles and triangles: excitation rates P inferred from the observations according to equation (9). The circles correspond to seismic observations from the BiSON network as quoted by Chaplin *et al.* (1998) and the triangles to that from the GOLF instrument as quoted by Baudin *et al.* (2005). Continuous and dashed lines correspond to predicted excitation rates computed as explained in section 5. The continuous (dashed resp.) line corresponds to a calculation in which χ_k is assumed Lorentzian (Gaussian resp.).

then probe a different region of the atmosphere. The BiSON instruments use the KI line whose height of formation is estimated at the height $h \approx 280$ km while the GOLF instrument uses the Na I D1 and D2 lines whose height of formation is estimated at the height $h \approx 340$ km (see Baudin *et al.* 2005). In Fig. 1, P_{obs} are calculated with mode masses evaluated at $h = 280$ km for the BiSON data set reduced by Chaplin *et al.* (1998) and at $h = 340$ km for the GOLF data set reduced by Baudin *et al.* (2005).

5. p mode excitation in the sun

The rate P at which energy is injected into solar p modes, was computed in Paper III and IV as follows: the calculations of ξ_r and ω_0 were performed on the base of a 1D solar model built according to Gough's (1977) non-local formulation of the mixing-length theory (GMLT hereafter). Values of ρ_0 and α_s were also obtained from this 1D GMLT's model. On the other hand, the quantities w , $\tilde{s}$ and Φ – that are related to the averaged properties of turbulent convection – were calculated directly from a 3D simulation of the solar surface with a spatial mesh grid of $253 \times 253 \times 163$ (see details in Paper III and IV). Furthermore the analytical models adopted for $E(k)$ and χ_k were also constraints by the 3D simulation.

It was shown from the 3D simulation that a Gaussian – usually used for modeling χ_k – is inadequate (Paper IV): a Lorentzian fits best the frequency dependence of χ_k .

As shown in Fig. 1 with the dashed line, the use of a Gaussian for χ_k underestimates P_{obs} while a Lorentzian results in values of P closer to the different solar seismic constraints.

This result then shows that, provided that such a non-Gaussian model is assumed, the present model of stochastic excitation is – for the sun – rather satisfactory, without adjustment of free parameters.

Some discrepancies still exist between P and P_{obs} as well as between the different seismic data analysis. In Fig. 1, P obtained with a Lorentzian is closer to P_{obs} inferred from the seismic data reduction by Baudin *et al.* (2005) whereas large discrepancies are obtained at high frequency with the values of P_{obs} inferred from Chaplin *et al.*'s (1998) data set. Furthermore we find that the height where χ_k in equation (9) is evaluated changes P_{obs} significantly (not shown). We then conclude that inferring excitation rates from the seismic observations remains difficult and provides constraints on P with still important uncertainties.

6. p mode excitation across the HR diagram

6.1 Main sequence stars

We consider a set of 3D simulations of solar-like oscillating stars. We compute internal structure of 1D stellar models and associated eigenfunctions consistent with the 3D simulations. We then compute, on the basis of the current theoretical model of stochastic excitation, the rates at which energy is expected to be injected into the modes. The calculation is performed as explained in Samadi *et al.* (2005a) and as summarized below: the calculation of ξ_r and ω_0 is performed using a 1D stellar model consistent with the simulations of stars. The total kinetic energy contained in the turbulent kinetic spectrum, $E(k)$, its depth dependence, and its k -dependence are obtained directly from the 3D simulation. For the eddy time-correlation function χ_k , we investigate both a Gaussian and a Lorentzian. Results are shown in Fig. 2 (left).

As mentioned in the introduction, the energy injected into the mode may be directly calculated from a 3D simulation of the outer layer of the star. This calculation has been done for the sun by Stein & Nordlund (2001) according to the expression proposed by Nordlund & Stein (2001) (their equation (74)), which corresponds to the direct calculation of $p dV$ work of the non-adiabatic gas and turbulent pressure (entropy and Reynolds stress) fluctuations on the modes. This method has been applied in Paper V to seven 3D simulations of the outer layers of main sequence stars. The result of this calculation is presented in Fig. 2 (left) by the triangles labeled as P_{3D} . These calculations are then compared with those obtained on the basis of the theoretical model of stochastic excitation.

A rather good agreement is found between P_{3D} and the calculations in which a Gaussian or an exponential (not shown) is assumed while a Lorentzian results in P much larger than P_{3D} .

In contrast to what was found in the case of the solar 3D simulation investigated in Paper IV (see section 5), the inferred ν dependencies of χ_k are far from a Lorentzian. This discrepancy is explained by the fact that the simulations investigated in Paper V have all a too small resolution compared to that considered in Paper IV. However, by investigating solar simulations with different resolutions, we find that, as the spatial resolution increases, χ_k tends towards a Lorentzian ν -dependence.

We find that P_{max} scales as $(L/M)^s$ where s is the slope of the scaling law, L is the luminosity and M is the mass of the 1D models associated with the 3D simulations. This result is not surprising: indeed, even though the ratio L/M is the ratio of two

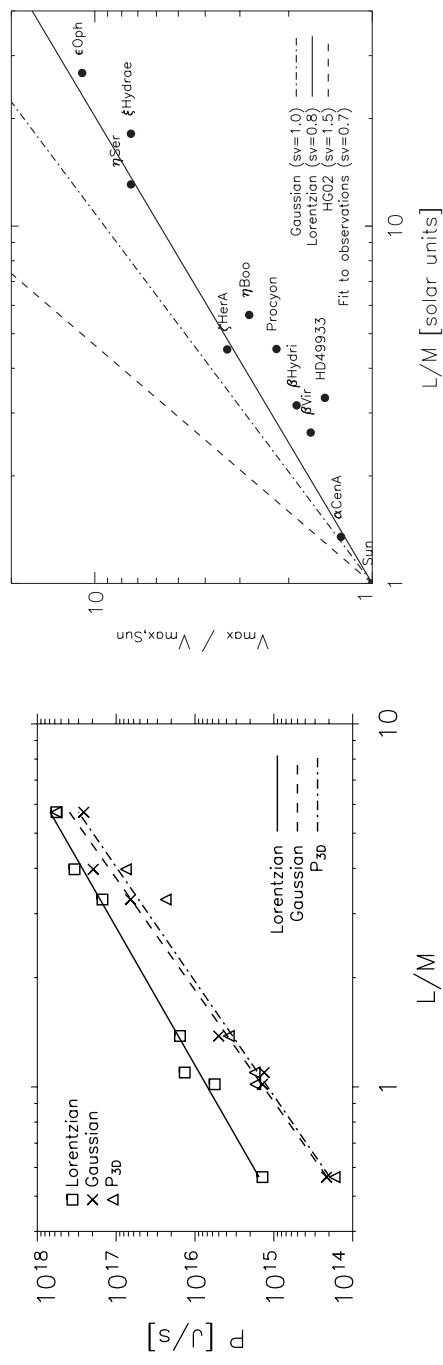


Figure 2. **Left:** $P_{\max}$ versus L/M where L is the luminosity and M is the mass of the 1D models associated with the 3D simulations. The triangles correspond to the direct calculations (labeled here as P_{30}) and the other symbols correspond to calculations based on the theoretical model of stochastic excitation and using the two forms of χ_k : the crosses assume a Gaussian and the squares a Lorentzian, respectively. The lines are the results of fitting each set of symbols with a power law of the form $(L/M)^s$ where s is the slope of the power law. **Right:** Same as left panel for $V_{\max}/V_{\max,\odot}$, the maximum of the mode amplitudes relative to the observed solar value ($V_{\max,\odot} = 27 \text{ cm s}^{-1}$). The filled circles correspond to the few stars for which solar-like oscillations have been detected in Doppler velocity. The solid and the dot dashed lines are the result of fitting the calculations performed with a Lorentzian and a Gaussian, respectively, with a power law of the form $(L/M)^{sv}$ where sv is the slope of the power law (see section 7.1). The dot dot dashed line is the result of fitting the observations with such power law. For comparison the dashed line shows the power law obtained by Houdek & Gough (2002).

global stellar quantities, it nevertheless characterizes essentially the stellar surface layers where the mode excitation is located since $L/M \propto T_{\text{eff}}^4/g$.

The value found for s is sensitive to the choice for χ_k : indeed the Gaussian results in $s = 3.3$ and the Lorentzian in $s = 2.6$. Seismic observations in the sun (see section 5) and in α Cen A tend to favor the Lorentzian over the Gaussian. For other stars for which stochastically excited p modes are expected, the Lorentzian is likely to remain the best form for χ_k (see section 7.1), but this needs to be confirmed with data providing constraints on the excitation rates P . The high-quality data of the COROT instrument will enable us to derive P_{max} as a function of L and M for the different target stars. These data will very likely enable us to determine the appropriate model for the eddy time-correlations (χ_k).

6.2 Pre-main sequence stars

p mode excitation has so far been studied in main sequence stars. We consider here a set of 1D models of pre-main sequence stars (PMS).

Our evolutionary models are computed with the CESAM code written by Morel (1997). The assumed physics is standard and very similar to that described in Lochard *et al.* (2005): the external boundary condition is defined in a simplified model atmosphere involving the Eddington $T(\tau)$ law. The adopted mean chemical composition for the sun is taken as $X_0 = 0.715$, $Y_0 = 0.267$, and $Z_0 = 0.018$. The convective flux is computed according to the classical prescription of the Mixing Length Theory (Böhm-Vitense (1958)). The mixing length value l_{MLT} is equal to $1.62 H_p$. No diffusion and no rotation are included in stellar model computations.

We have computed nine models of PMS stars with masses ranging from $M = 1 M_\odot$ to $M = 1.8 M_\odot$. Their locations in the HR diagram are shown in Fig. 3 (left). They are labeled from 1 to 9. The evolutionary tracks associated with those computed PMS models are also presented. The evolution of a PMS star may be decomposed into two different phases. The first phase corresponds to a fast contraction. During this phase the contraction time scale is of the same order as the dynamic time scale. The second phase, on the other hand, is characterized by a contraction time scale longer than the dynamic time scale. During this phase the usual quasi-static approximation then holds (see Suran *et al.* 2001) and the models computed by CESAM may be considered representative (except effects of fast rotation that are not included here).

Most of the models are chosen at a late stage of the second phase, before the stars reach the MS except model number 3 that is located at the beginning of the second phase.

Among our set of models, two of them (models 8 and 9) have $M \gtrsim 1.7 M_\odot$ and are thus representative of Herbig Ae stars. Furthermore both lie on the instability strip predicted by Dupret *et al.* (2004). All other models have a mass $M \lesssim 1.7 M_\odot$ and thus correspond to T-Tauri stars.

We next compute P_{max} for each model as follows: as for the other case (sections 5 and 6.1), the calculation of ξ_r and ω_0 is based on the 1D model and makes use of an adiabatic pulsation code. However the input quantities related to the spatial and time averages of the medium are obtained from the 1D models, which rely on the classical mixing-length. For the k -dependency of the kinetic energy spectrum, E , we assume as in Paper III and IV, the spectrum that fits best the spectrum obtained in the solar 3D simulation investigated in those papers. Finally we adopt a Lorentzian for χ_k .

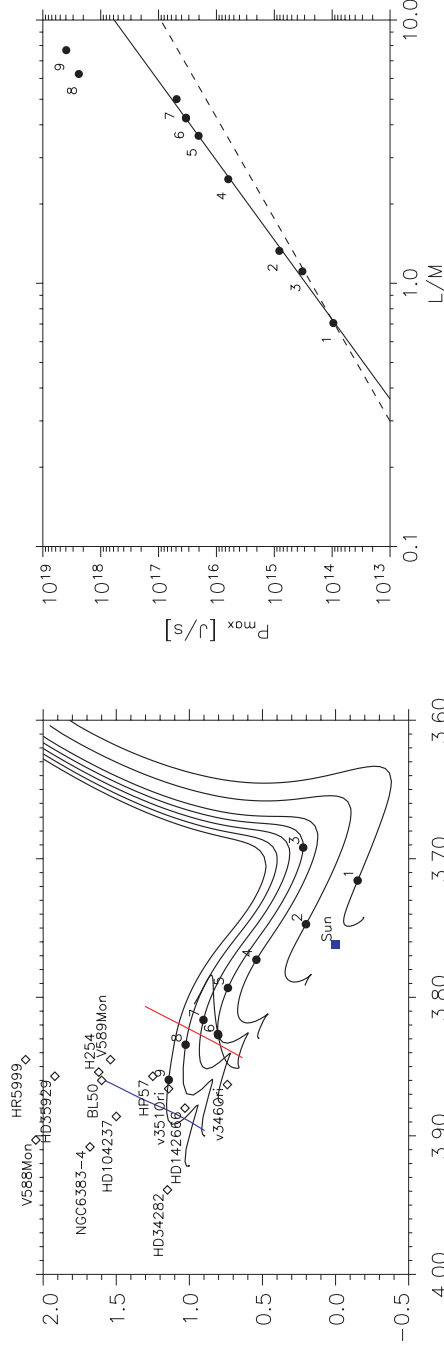


Figure 3. **Left:** Location of PMS stars on the HR diagram. The diamonds correspond to the PMS stars for which unstable modes have been detected (according to Table 2 of Marconi & Palla 2003). The filled circles with associated numbers correspond to computed PMS models with different ages and masses. The diagonal lines correspond to the red and blue edges of the instability strip as predicted by Dupret *et al.* (2004) for the radial modes. The evolutionary tracks of PMS models with masses ranging from $M = 1 M_{\odot}$ to $M = 1.8 M_{\odot}$ are also represented. **Right:** The filled circles correspond to the maximum of mode excitation rates obtained as explained in section 6.2 for the PMS models shown in the left panel. The continuous line is the result of fitting the set of dots with a power law of the form $(L/M)^s$ where s is the slope of the power law (models with number 8 and 9 have been discarded from the fit, see text). The dashed line corresponds to the power law $(L/M)^{2.6}$ obtained with calculations performed on the base of 3D simulation constraints and assuming a Lorentzian for χ_k (see section 6.1).

The results of the calculation of $P_{\max}$ are shown in Fig. 3 (right). Except for models 8 and 9, $P_{\max}$ may be nicely fitted with a power law of the form $(L/M)^s$ with $s = 3.3$. Hence, as for the main sequence stars, $P_{\max}$ increases with the ratio L/M according to a power law. This result is not surprising. Indeed, as for MS stars, p mode excitation occurs at the surface of the star that is characterized by a given T_{eff} and a given g (we recall that $L/M \propto T_{\text{eff}}^4/g$).

This slope of the power law is however somewhat larger than the one found with the set of main-sequence stars investigated in section 6.1. The difference is likely to be connected with the fact that the calculations in section 6.1 used constraints from the 3D simulations whereas those performed here rely on the mixing-length theory. However the physical reasons are not yet identified.

Models 8 and 9 result in larger $P_{\max}$ than one would expect with the fitted power law. In contrast with the other models, those two models have two separated super-adiabatic layers. The mode driving hence occurs within two different regions and is thus stronger than if those stars had only one super-adiabatic layer.

7. p mode amplitudes across the HR diagram

Calculation of the rms value of the mode surface velocity, v_s , requires the knowledge of the mode excitation rate (P) and the mode damping rate (η). Indeed according to equation (9), v_s^2 is related to P and η as:

$$v_s^2 = \frac{1}{\mathcal{M}} \frac{P}{\eta}, \quad (10)$$

where $\mathcal{M}$ is the mode mass evaluated at the photosphere (*i.e.*, at $r = R_{\text{eff}}$).

7.1 Main sequence stars

In Samadi *et al.* (2001b), calculations of the damping rates η based on Gough's time-dependent and non-local treatment of convection were performed for different stellar models with different values of L and M . On the other hand, as summarized in sections 6.1 and 6.2, $P_{\max}$ scales as $(L/M)^s$ where the slope s depends on the choice for χ_k . Then, according to equation (10), we may determine $V_{\max}(L, M)$ for the different power laws of $P_{\max}$. Each set of calculations is next fitted by a power law of the form $(L/M)^{sv}$. This was done in Paper V and the result is shown in Fig. 2 (right).

As for $P_{\max}$, the slope of the power law for $V_{\max}(L, M)$ depends on the assumptions for χ_k , that is $sv = 1.0$ when χ_k is Gaussian and $sv = 0.8$ when it is Lorentzian.

Solar-like oscillations have been detected in Doppler velocity in about ~ 10 stars. The maximum amplitudes of those oscillations are represented in Fig. 2 (right) with the filled circles. As already stressed by Kjeldsen & Bedding (1995), the amplitudes clearly scale approximately as $(L/M)^{sv}$.

“The observed dependence of $v_{\max}$ on the ratio L/M may be qualitatively explained as follows: As already noticed in Sect. 6.1, the maximum of the excitation rate, $P_{\max}$, scales as $(L/M)^s$. We recall that $L/M \propto T_{\text{eff}}^4/g$. The p mode excitation occurs at the surface of the star that is characterized by a given T_{eff} and a given g . Hence the larger T_{eff} , the larger the convective flux. Moreover the smaller g or the larger T_{eff} , the larger the pressure scale height and hence the larger the eddy size (see also

Kjeldsen & Bedding 1995). Now, the larger the eddy size or the larger the convective flux, the larger the excitation (c.f. Eq. 65 of Paper I)''.

A fit to the observations results in $sv \simeq 0.7$. As seen in Fig. 2 (right) calculations based on a Lorentzian for χ_k result in power law closer to that derived from the observations than those based on Gaussian. As a conclusion, as in the sun (see section 5) and in α Cen A, the observations tend to confirm that the Lorentzian is an approximately correct form for χ_k .

For comparison with previous results, the dashed line in Fig. 2 (right) shows the power law proposed by Houdek & Gough (2002). This power law clearly overestimates by a large amount the oscillation amplitudes observed in the stars more luminous than the sun. From this result, we conclude that the problem of the overestimation of the amplitudes of the solar-like oscillating stars more luminous than the sun is directly related to the choice of the model for χ_k .

7.2 Pre-main sequence stars

To compute the mode amplitudes associated with our PMS models we must in principle compute the mode damping rates. However for the high frequency p modes, the mode damping occurs mainly at the stellar surface. So we may expect that modes are damped at the same level as for the MS stars because MS and PMS stars have similar envelope properties. We may then expect that the mode amplitudes will vary as for the MS stars as $(L/M)^{sv}$, with a similar slope sv as found in section 7.1 (namely $sv = 0.7$).

Accordingly, $V_{\max}$ is expected to range between ~ 20 cm/s and ~ 120 cm/s for our set of stellar PMS models. To derive the mode amplitudes in terms of luminosity fluctuations, $(dL/L)_{\max}$, we may – as in Lochard *et al.* (2005) – apply a crude conversion based on the adiabatic approximations (see Kjeldsen & Bedding 1995). This puts $(dL/L)_{\max}$ in the range between ~ 2 ppm and ~ 15 ppm.

HARPS is currently among the best ground-based instruments devoted to the detection of solar-like oscillations. Indeed this instrument is able to detect solar-like oscillations in stars brighter than $m_v \sim 5$ (see e.g., Mosser *et al.* 2005). Unfortunately PMS stars are in general faint objects. For instance, the brightest known object has an apparent magnitude $m_v \sim 9$ (Marconi & Palla 2003). HARPS performances are still then too low to detect solar-like oscillations in the known PMS stars.

The CoRoT mission (see Baglin 2003) will be able to detect stochastically excited modes as low as in the sun (that is $(dL/L)_{\max} \simeq 4$ ppm), or in stars of apparent magnitude around $m_v \sim 6$ and with a signal-to-noise ratio (S/B) in the *power spectrum* better than 15. The instrument will be limited by the photon noise up to $m_v \sim 9$. For a star with magnitude $m_v \sim 8$, the (white) noise level will be of order 1.8 ppm, which will allow the detection of amplitudes larger than ~ 7 ppm with $S/B \gtrsim 15$. Above $m_v \sim 9$, the noise level is expected to increase very rapidly at a level difficult to predict now. Hence, the detection of solar-like oscillations in PMS stars is not guaranteed with the CoRoT instrument. On the other hand, a space based instrument with performances similar to that of the Eddington instrument (see Favata *et al.* 2000) will be able to detect solar-like oscillations in stars as faint as most of the known PMS stars.

8. Discussion

Although the current model of stochastic excitation results in excitation rates in good agreement with the values inferred from the solar seismic observations, there are still

important discrepancies with those derived from the recent seismic observations of α Cen A. In all cases the observations favor a Lorentzian for the eddy time-correlation function χ_k rather than a Gaussian. It is not yet clear whether the remaining discrepancies between the observations and the theory are related to the way the constraints are inferred from the observations or the assumptions and the simplifications adopted in the current theory.

Among the different simplifications involved in the theory, we believe that two of them are the most significant: First the model presented here is based on the strong assumption that the excitation is spatially isotropic (the medium anisotropy is only taken into account in average). Second, the calculation is based on the adiabatic (real) mode eigenvectors. Non-adiabatic effects are important in the region where the excitation occurs. They are expected to change the behavior of the eigenvectors in that region (work in progress).

Acknowledgements

RS and MJG acknowledge support by Programme National de Physique Stellaire (PNPS, France) and by the Aryabhata Research Institute of Observational Sciences (ARIES, Nainital, India). We thank C. Catala for fruitful discussions and Y. Lebreton for valuable help in building the PMS models.

References

- Baglin, A. 2003, *Advances in Space Research*, **131**(2), 345.
 Balmforth, N. J. 1992, *MNRAS*, **255**, 639.
 Baudin, F., Samadi, R., Goupil, M. *et al.* 2005, *A&A*, **433**, 349.
 Bedding, T. R., Kjeldsen, H., Butler, R. P. *et al.* 2004, *ApJ*, **614**, 380.
 Böhm-Vitense, E. 1958, *Zeitschr. Astrophys.*, **46**, 108.
 Butler, R. P., Bedding, T. R., Kjeldsen, H. *et al.* 2004, *ApJ*, **600**, L75.
 Chaplin, W. J., Elsworth, Y., Isaak, G. R. *et al.* 1997, *MNRAS*, **288**, 623.
 Chaplin, W. J., Elsworth, Y., Isaak, G. R. *et al.* 1998, *MNRAS*, **298**, L7.
 Christensen-Dalsgaard, J., Frandsen, S. 1983, *Solar Physics*, **82**, 469.
 Dupret, M.-A., Grigahcène, A., Garrido, R., Gabriel, M., Scuflaire, R. 2004, *A&A*, **414**, L17.
 Favata, F., Roxburgh, I., Christensen-Dalsgaard, J. 2000, In *The Third MONS Workshop: Science Preparation and Target Selection*, p. 49–54.
 Goldreich, P., Keeley, D. A. 1977, *ApJ*, **212**, 243.
 Goldreich, P., Murray, N., Kumar, P. 1994, *ApJ*, **424**, 466.
 Gough, D. O. 1977, *ApJ*, **214**, 196.
 Grigahcène, A., Dupret, M.-A., Gabriel, M., Garrido, R., Scuflaire, R. 2005, *A&A* (in press).
 Houdek, G., Balmforth, N. J., Christensen-Dalsgaard, J., Gough, D. O. 1999, *A&A*, **351**, 582.
 Houdek, G., Gough, D. O. 2002, *MNRAS*, **336**, L65.
 Kjeldsen, H., Bedding, T. R. 1995, *A&A*, **293**, 87.
 Kjeldsen, H., Bedding, T. R. 2001, In: *Proceedings of the SOHO 10/GONG 2000 Workshop: "Helio- and Asteroseismology at the dawn of the millennium"*, Vol. ESA SP-464, 361.
 Kjeldsen, H., Bedding, T. R. 2004, In: *Proceedings of the SOHO 14/GONG 2004 Workshop: "Helio- and Asteroseismology: Towards a Golden Future"*, Vol. ESA SP-559, 101.
 Libbrecht, K. G. 1988, *ApJ*, **334**, 510.
 Lochard, J., Samadi, R., Goupil, M.-J. 2005, *A&A* (in press).
 Marconi, M., Palla, F. 2003, *AP&SS*, **284**, 245.
 Morel, P. 1997, *A&AS*, **124**, 597.
 Mosser, B., Bouchy, F., Catala, C. *et al.* 2005, *A&A*, **431**, L13.
 Nordlund, Å., Stein, R. F. 2001, *ApJ*, **546**, 576.

- Samadi, R. 2001, In *SF2A-2001: Semaine de l'Astrophysique Francaise, E148* (astro-ph/0108363).
- Samadi, R., Goupil, M. 2001, *A&A*, **370**, 136 (Paper I).
- Samadi, R., Goupil, M., Lebreton, Y. 2001a, *A&A*, **370**, 147 (Paper II).
- Samadi, R., Houdek, G., Goupil, M.-J., Lebreton, Y., Baglin, A. 2001b, In: *1st Eddington Workshop: Stellar Structure and Habitable Planet Finding*, Vol. ESA SP-485 (astro-ph/0109174), 87–94.
- Samadi, R., Goupil, M.-J., Houdek, G. 2002, *A&A*, **395**, 563.
- Samadi, R., Nordlund, Å., Stein, R. F., Goupil, M. J., Roxburgh, I. 2003a, *A&A*, **403**, 303. (Paper III).
- Samadi, R., Nordlund, Å., Stein, R. F., Goupil, M. J., Roxburgh, I. 2003b, *A&A*, **404**, 1129 (Paper IV).
- Samadi, R., Goupil, M. J., Baudin, F. et al. 2004, In: *Proceedings of the SOHO 14/GONG 2004 Workshop: "Helio- and Asteroseismology: Towards a Golden Future"*, Vol. ESA SP-559, 615.
- Samadi, R., Georgobiani, D., Trampedach, R. et al. 2005a, submitted to *A&A* (Paper V).
- Samadi, R., Kupka, F., Goupil, M.-J., Lebreton, Y., van't Veer-Menneret, C. 2005b, submitted to *A&A*.
- Stein, R., Georgobiani, D., Trampedach, R., Ludwig, H.-G., Nordlund, Å. 2004, *Solar Physics*, **220**, 229.
- Stein, R. F., Nordlund, Å. 2001, *ApJ*, **546**, 585.
- Suran, M., Goupil, M., Baglin, A., Lebreton, Y., Catala, C. 2001, *A&A*, **372**, 233.
- Xiong, D. R., Deng, L., Cheng, Q. L. 1998, *ApJ*, **499**, 355.

HD 101065, the Most Peculiar Star: First Results from Precise Radial Velocity Study

D. E. Mkrtichian¹ & A. P. Hatzes²

¹ARCSEC, Sejong University, Seoul 143 747, Korea and Astronomical Observatory, Odessa National University, Shevchenko Park, Odessa 65014, Ukraine.

e-mail: david@arcsec.sejong.ac.kr

²Thüringer Landessternwarte Tautenburg, Sternwarte 5, D-07778, Tautenburg, Germany.

Abstract. In this paper we discuss the prospects for asteroseismology with spatial resolution and motivate studies of the most chemically peculiar roAp star HD 101065. We present the first results from a high-precision radial velocity (RV) study of HD 101065 based on data spanning four nights that were acquired using the HARPS echelle-spectrometer at the ESO 3.6 m telescope. The analysis of individual nights showed the amplitude and phase modulation of the dominant mode. The analysis of the whole data set showed the presence of multi-periodic oscillations with two groups of equally-spaced modes. We find $\Delta\nu = 65.2 \mu\text{Hz}$ and $\delta\nu = 7.3 \mu\text{Hz}$ for the large and the small spacing, respectively. HD 101065 is the only roAp star to show the existence of two groups of $l = 0, 2$ and $l = 1, 3$ excited modes.

Key words. Stars: oscillations, roAp stars.

1. Prospects for asteroseismology with spatial resolution of roAp stars

The main difference in observations in asteroseismology and helioseismology is that we cannot resolve the disks of most stars. As a consequence, cancellation effects reduce the integrated light and velocity amplitudes of pulsations causing only the lowest degree non-radial pulsation (NRP) modes ($l = 1\text{--}4$) to be detected in slowly rotating stars. In rapidly rotating stars higher degree modes ($l \approx 4\text{--}20$) can be detected by the spectral distortions in the rotationally broadened spectral lines (see Vogt & Penrod 1983). By comparison, resolved disk observations of the sun allow the detection of pulsation modes up to $l \approx 600$. The narrow l -range of modes visible in stellar disk-integrated studies decreases the diagnostic potential of asteroseismology. Furthermore, mode identification methods based on disk-integrated photometric or spectroscopic measurements do not work well. This is the most severe problem in asteroseismology.

The roAp stars are a unique group of pulsators that present us with advantages for asteroseismic studies compared to other stars. There are several similarities with helioseismology. First, the excited spectrum of roAp stars are due to high-overtone p -modes showing equal-spacing of pulsation frequencies (Kurtz 1990). Second, the identification of high-overtone p -modes is relatively simple due to the equal spacing of modes that makes the discrimination of $l = \text{odd}$ and $l = \text{even}$ groups of modes

possible. The oscillations in roAp are easier to detect as they have higher amplitudes which are relatively stable compared to oscillations in the sun and the solar-type stars.

Another important advantage in asteroseismology of roAp stars is the possibility of the spatial filtration of NRP modes (Mkrtichian 1994). Spatial filtration exploits the fact that elements on the surface of Ap (including roAp) stars are distributed inhomogeneously. Spectroscopic radial velocity measurements of elements that have a spotted distribution on the surface measure a pulsational velocity that can originate from a small region on the star. The cancellation effects are thus smaller and higher degree modes can be detected as well as low amplitude ones. Recent work has demonstrated the efficiency of radial velocity spectroscopy for detection of low-amplitude modes (Hatzes & Mkrtichian 2004; Mkrtichian & Hatzes 2005; Kurtz *et al.* 2004).

There is growing evidence of the presence of a standing wave with a radial node in the atmospheres of some roAp stars. Using RV measurements of α Cir over narrow wavelength intervals Baldry *et al.* (1999) found that about 15% of the wavelength bands pulsated 180° out-of-phase with respect to most other lines. A similar result was found in 33 Lib by Mkrtichian *et al.* (2003) based on RV measurements of individual spectral lines. In particular they showed that Nd II and Nd III lines pulsate exactly out-of-phase. The existence of additional monotonous phase changes across the atmosphere was also seen in 33 Lib by Mkrtichian *et al.* (2003) and Kurtz *et al.* (2004). The vertical scale of the acoustic wave was found to be comparable to the size of line-forming regions of the atmosphere. In the roAp stars with a vertical acoustic standing or running wave there should be a one-to-one mapping between the pulsation phase and amplitude of the mode with respect to atmospheric geometric depth. Information on the vertical structure can come from the pulsation phases and amplitudes of the individual spectral lines of different elements that are formed at different atmospheric depths, as well as from bisectors of strong lines or lines of stratified elements. This important fact gives a chance for accurate acoustic study of the atmosphere using acoustic cross-sections.

The spatial resolution throughout the atmosphere is determined by the accuracy of the pulsation phase and amplitude measurements. The longer the time base of the observing, the higher the accuracy of the phase and amplitude determination. Moreover, different modes are expected to have different depths for the acoustic node and different phase profiles across the atmosphere. These thus will probe different atmospheric layers.

Several years ago, we started a program “Asteroseismology with spatial resolution” in order to study the three-dimensional structure of the atmospheres of roAp stars. An important part of these studies is the verification of the diffusion theory for the abundance anomalies in Ap stars that can be done by the determination of the acoustic and chemical profiles of the atmosphere. The acoustic confirmation of vertical separation of elements will be the most robust confirmation of the diffusion theory when combined with other approaches.

2. HD 101065 project motivation

The star HD 101065 was discovered by Przybylski (1961) as an extremely chemically peculiar star with strong spectral lines of rare-earth elements. The four magnetic field measurements by Wolff & Hagen (1976) show a longitudinal magnetic field in the

range of $H_z = -2100$ to -2500 ± 450 (Gauss). No significant magnetic field variability was found with these data. Recent magnetic field measurements by Hubrig *et al.* (2004) yield a longitudinal magnetic field strength of half the value, or $H_z = -1014 \pm 72$ Gauss.

HD 101065 was found to be pulsating by Kurtz (1978) and is a prototype star of the roAp group. The pulsation spectrum of the star was studied in detail by Martinez & Kurtz (1990) who found multiperiodic oscillations with amplitude and phase modulations of the dominant mode; however, the equal spacing of frequencies was not reliably determined. The detailed historical review on studies of HD 101065 (also known as Przybylski's star) and the problems with the determination of its main stellar parameters were given by Kurtz & Martinez (2000).

The chemical abundance investigation of this star is intriguing. In early work, Wegner & Petford (1974) identified the spectral lines of 51 chemical elements. Cowley *et al.* (2000) performed abundance determination for 54 elements and found that lanthanoids were overabundant by 4–5 dex. Cowley *et al.* (2004) also confirmed with a high degree of confidence the presence of lines of Tc and the short-lived (half-life $\tau = 17.7$ years) element Pm. The abundance and the isotopic content of Li was determined by Shavrina *et al.* (2004). Gopka *et al.* (2004) estimated the abundance limits of Tc, Pb, Bi and increased the number of elements with known abundances up to 59. They also identified the lines of all radioactive elements with atomic numbers from $Z = 84$ to $Z = 99$, except for $Z = 85$ and 87 (At and Fr).

The presence of short-living radioactive elements in Przybylski's star is enigmatic. Cowley *et al.* (2004) proposed that some unrecognized processes, perhaps flare activities, were taking place in the atmospheres of Przybylski's star. On the other hand, Gopka *et al.* (2004) suggested that the existence of elements with $Z < 92$ is due to the natural radioactive decay of thorium and uranium stratified in the upper atmosphere. The existence of the elements with $Z > 92$ is due to neutron capture in the stratified layers of the atmosphere with large overabundance of Th and U. In spite of 43 years of investigations, the extraordinary chemical anomalies of Przybylski's star are still unexplained.

We started a detailed study of Przybylski's star with the goals:

- to find and identify the *p*-mode oscillation spectrum,
- to determine the acoustic cross-sections of the atmosphere in order to confirm the existence of an acoustic node in line-forming atmosphere,
- to check for the presence of vertical stratifications of elements and thus confirm diffusion theory.

If the diffusion theory is valid, then there should be vertical stratifications of elements and with the majority of peculiar heavy elements levitated in the upper atmosphere. Spectral lines of elements formed on opposite sides of an acoustic node should pulsate with opposite phase to lines of non-peculiar elements formed in the lower atmosphere, similar to what was found in Nd II and Nd III in 33 Lib (Mkrtychian *et al.* 2003). The acoustic check for stratification of peculiar elements in upper layers should be the most accurate confirmation of the diffusion theory. Here we report the first results from four consecutive nights of RV measurements for HD 101065.

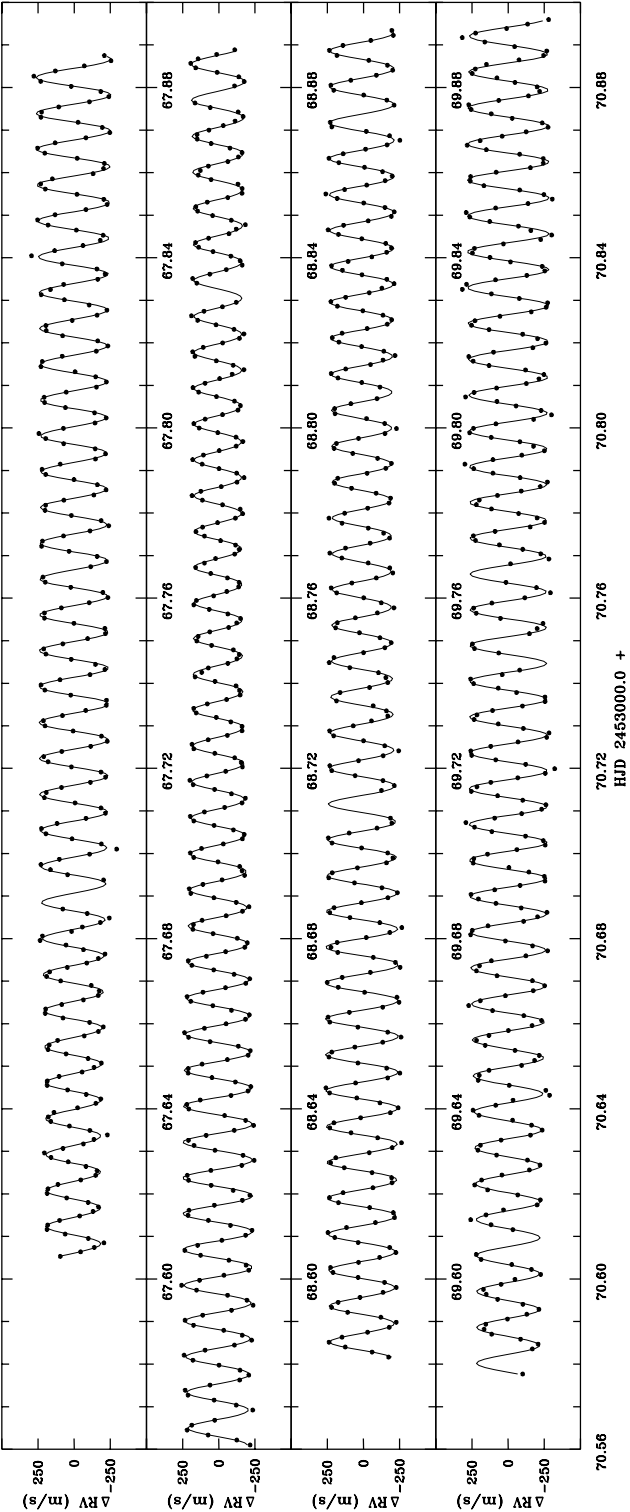


Figure 1. The “integrated” (over the interval 5000–6200 Å) radial velocities for HD 101065 on the four nights.

3. First results – p -mode spectrum of HD 101065

HD 101065 was observed on four consecutive nights (3rd–6th March 2004) using the HARPS echelle-spectrometer of the 3.6-m telescope of ESO. The resolving power $R = 115\,000$ in the wavelength range $\lambda\lambda 3800\text{--}6900$ and the iodine absorption cell was used. To achieve the maximum radial velocity (RV) precision all spectral lines in the full spectral region covered by iodine cell were used. The radial velocity data is shown in Fig. 1. The preliminary time-series analysis performed for individual nights confirm the dominant $f_1 = 1.3728$ mHz mode. In order to investigate the amplitude and phase variability of the dominant f_1 mode, each night was divided into time segments consisting of 8–10 pulsation cycles and the best fit to the data of each segment was then found. Figure 2 shows variations of the amplitude and phase. Both data are variable with $f_m = 1.36$ day modulation period that can be considered as the star's probable rotation period. To search for low-amplitude modes we removed the variable contribution of the dominant mode from the data of each night. The Discrete Fourier Transform (DFT) analysis of the merged residuals revealed 4 additional modes that are shown in Fig. 3. The frequencies and mean amplitudes of modes are given in Table 1.

4. Discussion and conclusion

A remarkable feature of the oscillation spectrum of HD 101065 is the existence of two groups of low (f_3 and f_4) and high-frequency (f_1 , f_2 and f_5) modes equally-spaced within each group by the mean value $\Delta\nu = 32.6\,\mu\text{Hz}$.

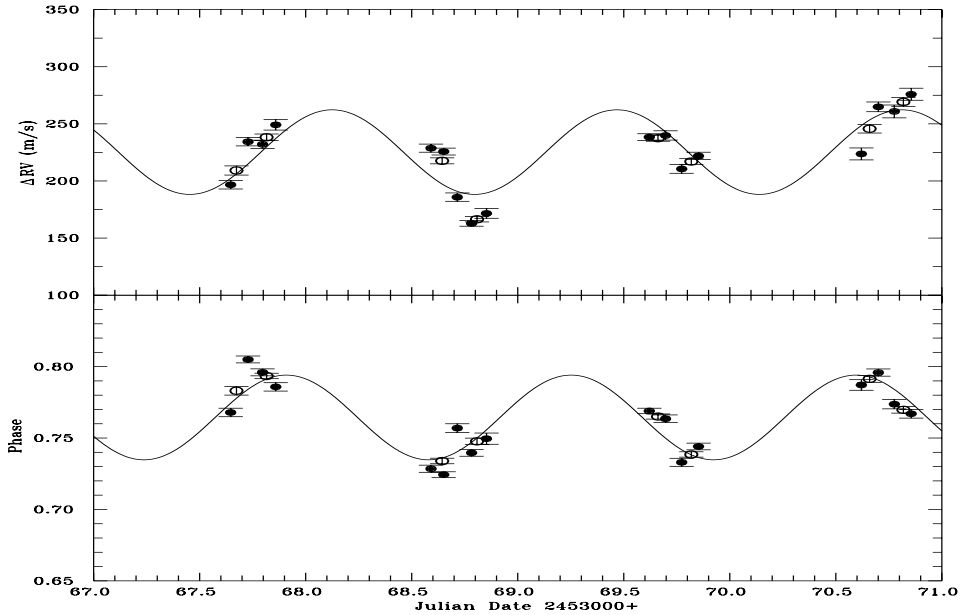


Figure 2. The modulation of amplitude (top panel) and phase (bottom panel) of the dominant $f_1 = 1372.8\,\mu\text{Hz}$ mode.

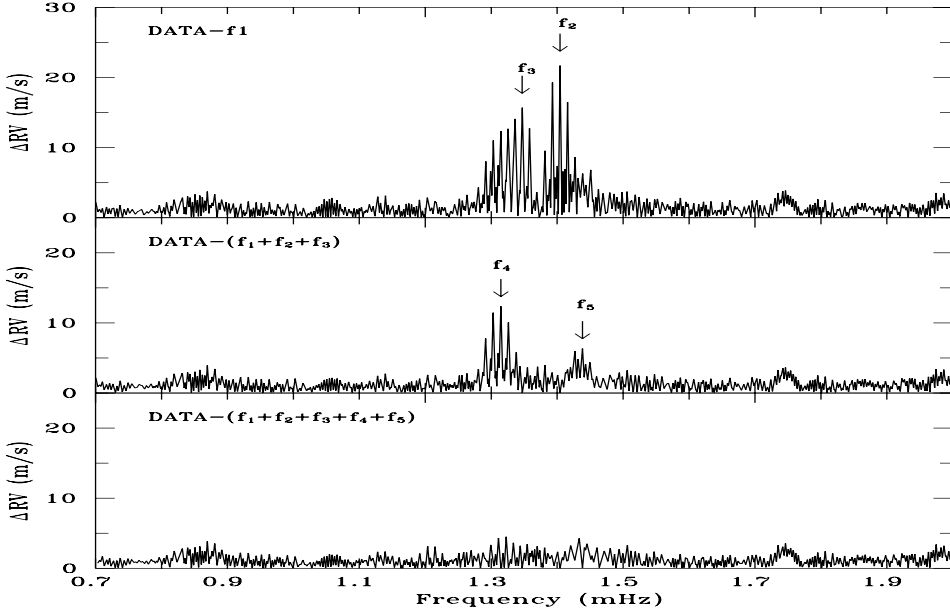


Figure 3. Consecutive prewhitening of the amplitude spectra. The top panel shows the DFT after removing the contribution of $f_1 = 1.3728$ mHz. Lower panels are the DFT after removal of the next dominant mode.

Table 1. Frequencies f_1 – f_6 detected in RV analysis.

No	Frequency μHz	K m s^{-1}	σ m s^{-1}
f_1	1372.8	221.3	1.92
f_2	1404.4	20.6	1.03
f_3	1314.6	12.58	0.89
f_4	1347.0	11.61	0.99
f_5	1438.4	6.43	0.96
$f_6 = 2 \times f_1$	2745.2	6.22	0.90

The echelle-diagram for p -mode spectrum plotted with a frequency spacing of $65.2 \mu\text{Hz}$ is shown in Fig. 4. The regularly spaced pattern of two groups of modes f_1, f_2, f_5 and f_3, f_4 that belong to modes of opposite l -parity is well visible. Modes with the same parity $l = 1, 3$ and $l = 0, 2$ are shifted in spacing on the echelle-diagram with respect to each other by the mean value $\delta\nu = 7.3 \mu\text{Hz}$.

This echelle-diagram has similarities with the p -mode spectrum of the sun and other solar-like stars (e.g., Eggenberger *et al.* 2004) in that it shows both large and small separation in the spacing of modes. We conclude, that two odd and two even groups of modes are excited in Przybylski's star. A comparison with echelle-diagrams for roAp stars (Mkrtichian & Hatzes 2000, 2005) shows that HD 101065 is currently the first roAp star for which the spectrum of $l = 0, 1, 2$ and 3 excited modes has been detected.

Detailed results on the study of p -mode spectrum of Przybylski's star is in preparation and will be published elsewhere.

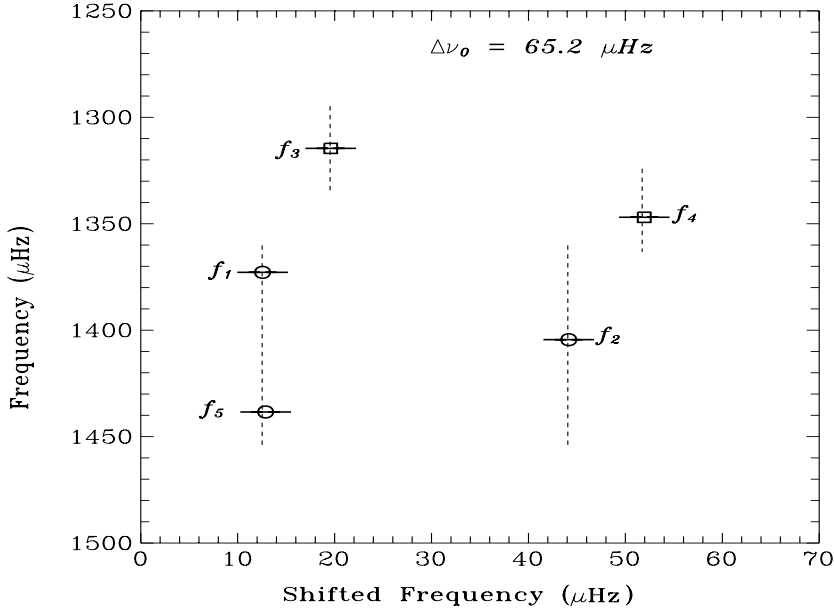


Figure 4. The echelle-diagram for HD 101065.

References

- Baldry, I. K., Viskum, M., Bedding, T. R., Kjeldsen, H., Frandsen, S. 1999, *MNRAS*, **302**, 381.
 Cowley, C. R., Ryabchikova, T., Kupka, F., Bord, D. J., Mathys, G., Bidelman, W. P. 2000, *MNRAS*, **317**, 299.
 Cowley, C. R., Bidelman, W. P., Hubrig, S., Mathys, G., Bord, D. J. 2004, *A&A*, **419**, 1087.
 Eggenberger, P., Charbonnel, C., Talon, S., Meynet, G., Maeder, A., Carrier, F., Bourban, G. 2004, *A&A*, **417**, 235.
 Gopka, V. F., Yushchenko, A. V., Shavrina, A. V., Mkrtichian, D. E., Hatzes, A. P., Andrievsky, S. M., Chernysheva, L. V. 2004, *The A-star Puzzle, Proc. IAU Symp.*, **224**, 734.
 Hatzes, A. P., Mkrtichian, D. E. 2004, *MNRAS*, **351**, 663.
 Kuchowicz, B. 1971, *Astrophys. Lett.*, **9**, 85.
 Kurtz, D. W. 1978, *IBVS*, **No 1436**.
 Kurtz, D. W. 1990, *ARA&A*, **28**, 607.
 Kurtz, D. W., Martinez, P. 2000, *Baltic Astronomy*, **9**, 253.
 Kurtz, D. W., Elkin, V. G., Mathys, G., Riley, J., Cunha, M. S., Shibahashi, H., Kambe, E. 2004, *The A-Star Puzzle, Proceedings of IAU Symp.*, **224**, 343.
 Martinez, P., Kurtz, D. W. 1990, *MNRAS*, **242**, 636.
 Mkrtichian, D. E. 1994, *Solar Physics*, **152**, 275.
 Mkrtichian, D. E., Hatzes, A. P. 2000, *PASPC*, **203**, 455.
 Mkrtichian, D. E., Hatzes, A. P., Kanaan, A. 2003, *MNRAS*, **345**, 781.
 Mkrtichian, D. E., Hatzes, A. P. 2005, *A&A*, **430**, 263.
 Przybylski, A. 1961, *Nature*, **189**, 739.
 Shavrina, A. V., Polosukhina, N., Khan, S., Pavlenko, Ya., Khalak, V., Wade, G. A., Quinet, P., Mikhailitska, N., Yushchenko, A., Gopka, V., Hatzes, A., Mkrtichian, D., Veles, A. 2004, *The A-star Puzzle, Proc. IAU Symp.*, **224**, 711.
 Vogt, S. S., Penrod, G. D. 1983, *ApJ*, **299**, 653.
 Wegner, G., Petford, A. D. 1974, *MNRAS*, **168**, 557.
 Wolff, S. C., Hagen, W. 1976, *PASP*, **88**, 119.

The Nainital–Cape Survey: A Search for Variability in Ap and Am Stars

Santosh Joshi

*Aryabhatta Research Institute of Observational Sciences (ARIES), Manora Peak,
Nainital 263 129, India.
e-mail: santosh@aries.ernet.in*

Abstract. The “Nainital–Cape Survey” program for searching photometric variability in chemically peculiar (CP) stars was initiated in 1997 at ARIES, Nainital. We present here the results obtained to date. The Am stars HD 98851, HD 102480, HD 13079 and HD 113878 were discovered to exhibit δ Scuti type variability. Photometric variability was also discovered in HD 13038, for which the type of peculiarity and variability is not fully explained. The null results of this survey are also presented and discussed.

Key words. Stars: Ap—stars: Am—stars: δ Scuti—individual: HD 98851—stars: individual: HD 102480—stars: individual: HD 13079—stars: individual: HD 13038—stars: individual: HD 113878.

1. Introduction

A collaborative survey program for searching variability in Ap (A-peculiar) and Am (A-metallic) stars between the Aryabhatta Research Institute of Observational Sciences, Nainital (formerly State Observatory Nainital), India and the South African Astronomical Observatory (SAAO) Cape Town, South Africa, was initiated in 1997 (Seetha *et al.* 2001). Prior to the Nainital–Cape Survey, a total of 31 rapidly oscillating Ap (roAp) stars were known, among which 28 belong to the southern sky. In order to investigate whether the difference in the observed number of roAp stars in both hemispheres corresponds to a reality or to an observational bias, the present survey was initiated to search for photometric variability in Ap and Am stars. The scientific goals of this survey were:

- to find a fresh perspective on the questions of variability in Ap and Am stars;
- to search for variability in stars lying in the same part of H–R diagram as that for roAp and δ -Scuti stars;
- to search for variability caused by g -modes, p -modes, or both; and
- to search for rapid oscillations in Am stars.

A number of other variability surveys are currently being performed, see, *e.g.*, Dhorokova *et al.* 2005, in this issue. To optimize the chances of discovering new variable stars, the strategy adopted in the Nainital–Cape Survey was to select candidates with Strömgren photometric indices similar to those of the known roAp stars (Martinez *et al.* 1991).

The Ap stars are characterized by anomalously strong lines of Si, Sr, Cr, Eu and rare earth elements in their optical spectra (North 1993). The effective temperature range of these stars varies from 7000 to 16,000 K. These stars possess relatively high magnetic fields (typically of the order of 1 to 10 kG). Ap stars have slow rotational velocities ($< 100 \text{ km sec}^{-1}$) as compared to normal A type stars which rotate with velocities in the range of 250 to 300 km sec^{-1} . The magnetic field strength, the strength of the spectral lines and the luminosity of Ap stars are all observed to vary periodically on time scales of days to several decades. This is explained with the help of the *oblique rotator model* (Stibbs 1950). The Ap stars showing photometric variability with periods in the range of $\sim 5\text{--}21$ min are known as rapidly oscillating Ap or roAp stars. In contrast to Ap stars, Am stars are non-magnetic stars with enhanced lines of metals in their spectra, as compared to normal stars of the same spectral classes (Titus & Morgan 1940; Roman *et al.* 1948; Smith 1974). Am stars are the coolest among chemically peculiar stars and have temperatures in a range of $\sim 6500\text{--}10,000$ K. The majority of (and possibly all) Am stars with spectral types later than A4 are members of spectroscopic binary systems with periods less than 100 days, whereas only a few normal A type stars are binaries with periods generally longer than 100 days. Some Am stars exhibit photometric variability similar to the δ Scuti stars, with periods from 18 min to as long as 7 hr.

This paper presents the results obtained from “Nainital–Cape Survey”. The paper is organized as follows. Section 2 summarizes the observations and the data analysis. The five variable stars discovered during the survey are presented in section 3, while the null results are presented and discussed in section 4. Conclusions are given in the last section.

2. Observations and data analyses

Most of the results presented here are based on the photometric observations. For some stars, we also carried out complementary spectroscopic observations. The photometric observations of the Ap and Am stars were carried out with a three-channel fast photometer (Ashoka *et al.* 2001; Sagar 1999) attached to the 104 cm Sampurnanand telescope at ARIES. Using this photometer, one can observe the programme star, a nearby comparison star and the sky-background measurement. As all the program stars for the survey are bright ($m_v < 10$ mag), it is very difficult to find a comparison star of similar magnitude and colour in the same field of view. Therefore, most of the time we decided to use only a single channel. The data reduction process comprised of four steps:

- (1) Visual inspection of the light curve to identify and remove all the bad data points.
- (2) Apply correction for the coincidence counting losses.
- (3) Measurement of sky brightness produced by interpolating the points with a piece-wise linear function.
- (4) Apply correction for the atmospheric extinction.

The time of the mid-points of the integrations were converted into Heliocentric Julian Dates (HJD) with an accuracy of 10^{-5} day (~ 1 sec). The reduced data thus comprises a time-series of HJDs and corrected magnitude with respect to the mean of that for the entire observations. The time-series data for all the objects were then analyzed using an algorithm (Kurtz 1985) based on Deeming’s DFT for equally spaced data (Deeming 1975).

The low resolution spectroscopic observations (in the spectral range of 3500–7400 Å) were also carried out from ARIES. We used for that purpose a $1\text{ k} \times 1\text{ k}$ CCD detector of $24\mu \times 24\mu$ pixel size mounted on the HR-320 spectrograph with a linear dispersion of $\sim 2.4\text{ Å}$ per pixel. The spectra were taken using a 300 lines/mm grating and a 3 mm circular aperture. The intermediate resolution spectroscopic observations were obtained from the Vainu Bappu telescope at Kavalur, using the Opto-Mechanics Research (OMR) spectrograph. In this case, a grating of 1200 lines/mm and 2.5 mm long slit was used for intermediate resolution spectroscopy, which yields a dispersion of 1.3 Å per pixel. Preliminary spectroscopic data analysis was performed using IRAF, developed by the National Optical Astronomy Observations (NOAO).

3. Results

3.1 HD 98851

HD 98851 is of the spectral type F1 with a visual magnitude of 7.4. The Strömgren photometric indices for this star are: $b - y = 0.199$, $m_1 = 0.222$, $c_1 = 0.766$ (ESA 1997). HD 98851 was observed for a total of 10 photometric nights. The two main periods of 75 min and 150 min were discovered by Joshi *et al.* (2000). Figure 1 shows one light curve of this star. The alternating high and low amplitude cycles indicate a nearly sub-harmonic period ratio of 2:1. Such a phenomenon was not known in any Am or δ Scuti stars and to our knowledge this was discovered for the first time. The complementary spectroscopic analyses show that HD 98851 is of spectral type F1IV (Joshi *et al.* 2003). A similar type of the light curve and identical periodicity was reported from Xinglong Station of the National Astronomical Observatory, China (NAOC), using a three-channel fast photometer attached to an 85 cm telescope (Zhou 2001).

3.2 HD 102480

HD 102480 has F1 spectral type and a visual magnitude of 8.40. The Strömgren photometric indices for this star are $b - y = 0.211$, $m_1 = 0.204$, $c_1 = 0.732$ (ESA 1997). Variability in HD 102480 was discovered and confirmed from ARIES by Joshi *et al.*

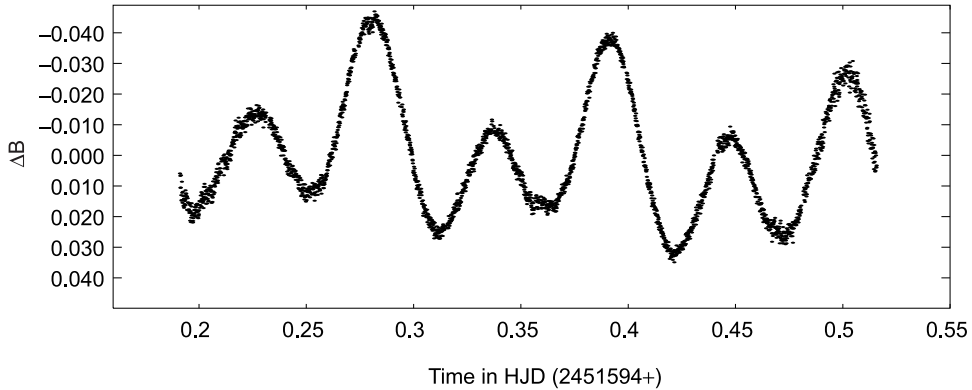


Figure 1. A light curve of HD 98851 obtained on HJD 2451594.

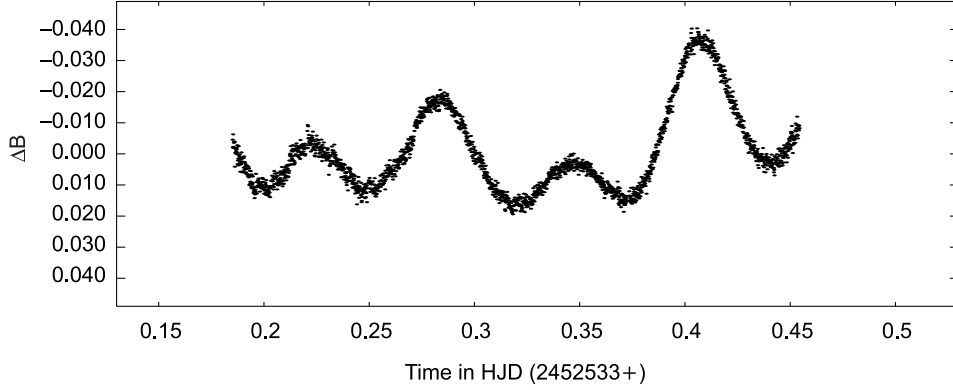


Figure 2. A typical light curve of HD 102480 obtained on HJD 2452533.

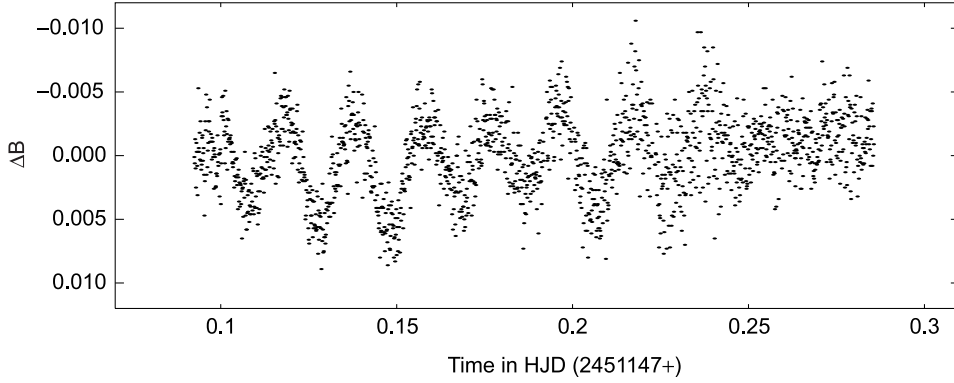


Figure 3. A typical light curve of HD 13038 obtained on HJD 2451147.

(2002). This star was observed for a total of 6 photometric nights. A typical light curve of this star is shown in Fig. 2. HD 102480 is pulsating with two main periods of 2.6 hr and 1.4 hr. The alternating high and low amplitude variations are similar to those of HD 98851. A nearly harmonic frequency ratio close to 2:1 is conspicuous. The complementary spectroscopic analyses shows that HD 102480 is of spectral type F3III/IV (Joshi *et al.* 2003).

3.3 HD 13038

HD 13038 is classified as a spectral type A3 star in the Henry Draper catalogue, having Strömgren colours $b - y = 0.105$, $m_1 = 0.220$, $c_1 = 0.848$ and $\beta = 2.86$ (ESA 1997). Variability in this star was discovered by Martinez *et al.* (1999a). HD 13038 was observed photometrically during 4 photometric nights from ARIES, see Fig. 3 for a typical light curve. This star is pulsating mainly with two close periods of 28 and 34 min. The amplitude modulation is caused by the beating of these periods, which lie between the typical pulsation period range of the roAp and δ Scuti stars. Recently, Elkin *et al.* 2005 have discovered a new luminous roAp star (HD 116114) with a pulsation period of 21 min, which is the longest period among all the known roAp stars.

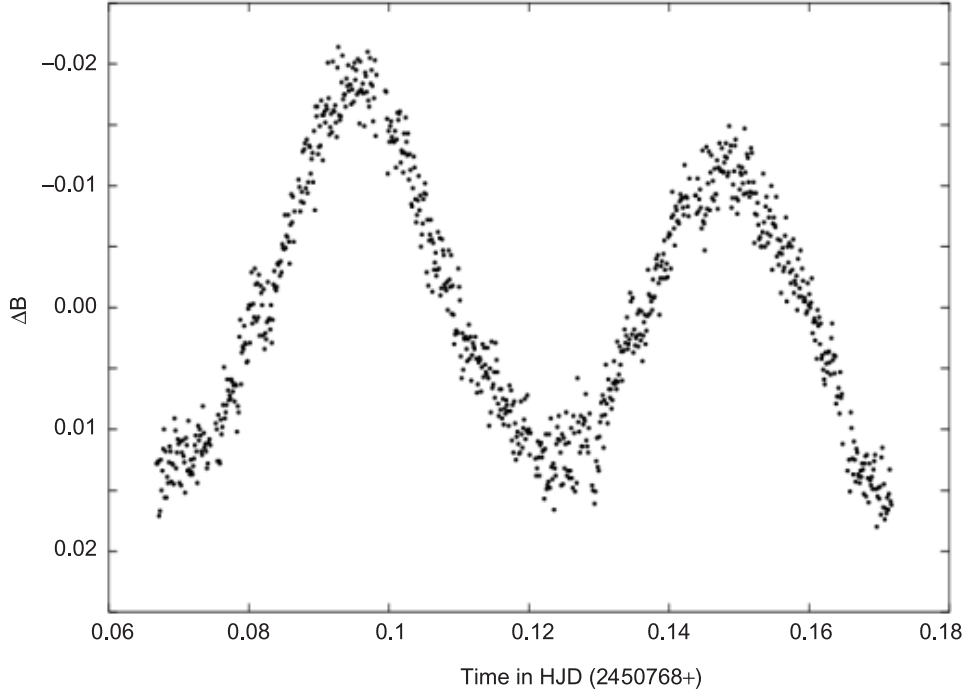


Figure 4. A typical light curve of HD 13079 obtained on HJD 2450768.

If HD 13038 is classified as an Ap star, then it will contribute in resolving a number of important questions related to the roAp and δ Scuti stars.

3.4 HD 13079

HD 13079 is a visual double star. The Hipparcos H_p magnitudes for the two components are 8.989 ± 0.007 and 11.31 ± 0.05 . The spectral type of HD 13079 is F0, and the Strömgren photometric indices for the combined light of both stars are $b - y = 0.203$, $m_1 = 0.211$, $c_1 = 0.672$ and $\beta = 2.759$ (ESA 1997). HD 13079 was observed photometrically from ARIES and the Vainu Bappu telescope (VBT), for a total of 5 photometric nights. Figure 4 shows a typical light curve of this star. The data analyses on HD 13079 show that this star is pulsating with a period of 78 min, which corresponds to a fundamental period of δ Scuti type variables. Spectroscopic analyses have shown that HD 13079 is an evolved Am star (Martinez *et al.* 1999b). The possible/probable change in the fundamental period will tell us more about the evolutionary state of this star, making this discovery very important from the asteroseismological point of view.

3.5 HD 113878

HD 113878 is a star of visual magnitude 7.8. The Strömgren photometric indices for this star are $b - y = 0.219$, $m_1 = 0.257$, $c_1 = 0.729$, $\beta = 2.45$ (ESA 1997). The spectral type of this star is F1/F3V/F3, on the basis of the CaII Kline, the H line and the metal lines respectively, suggesting that HD 113878 is a marginal Am star. HD 113878

was observed photometrically from ARIES, for a total of 8 photometric nights. The time-series analyses show that the star pulsates with a period of around 2.3 hr, which is again typical of δ Scuti stars, see Fig. 5 for a typical light curve. Abt (1984) has shown that the above stars HD 98851, HD 102480, HD 113878, as well as another one (HD 104202C) do not show the classical range of more than 0.5 spectral class between the K-line type and the metallic-line types. They are however classified as Am stars by their strong Sr II lines, weak $\lambda 4226$ CaI line, and other indications of their general abnormality. Among these stars, HD 98851, HD 102480 and HD 113878 show photometric variability. The star HD 104202C has to be observed for variability. If this star exhibits variability patterns similar to the other three objects, then these variable stars

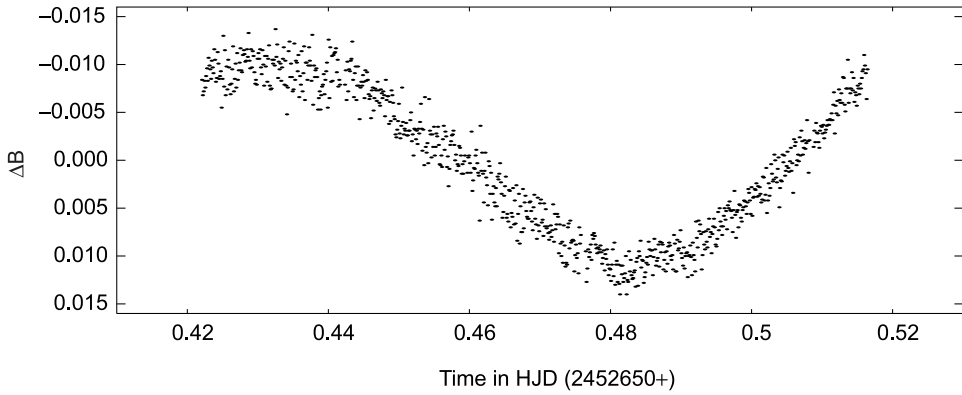


Figure 5. A typical light curve of HD 113878 obtained on HJD 2452650.

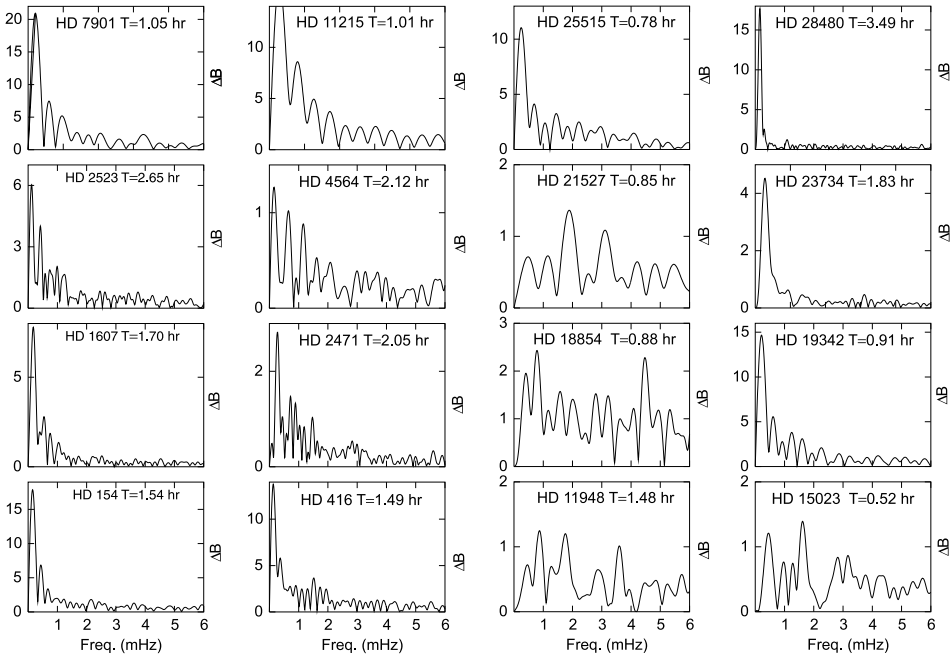


Figure 6. Negative results from "Nainital-Cape Survey".

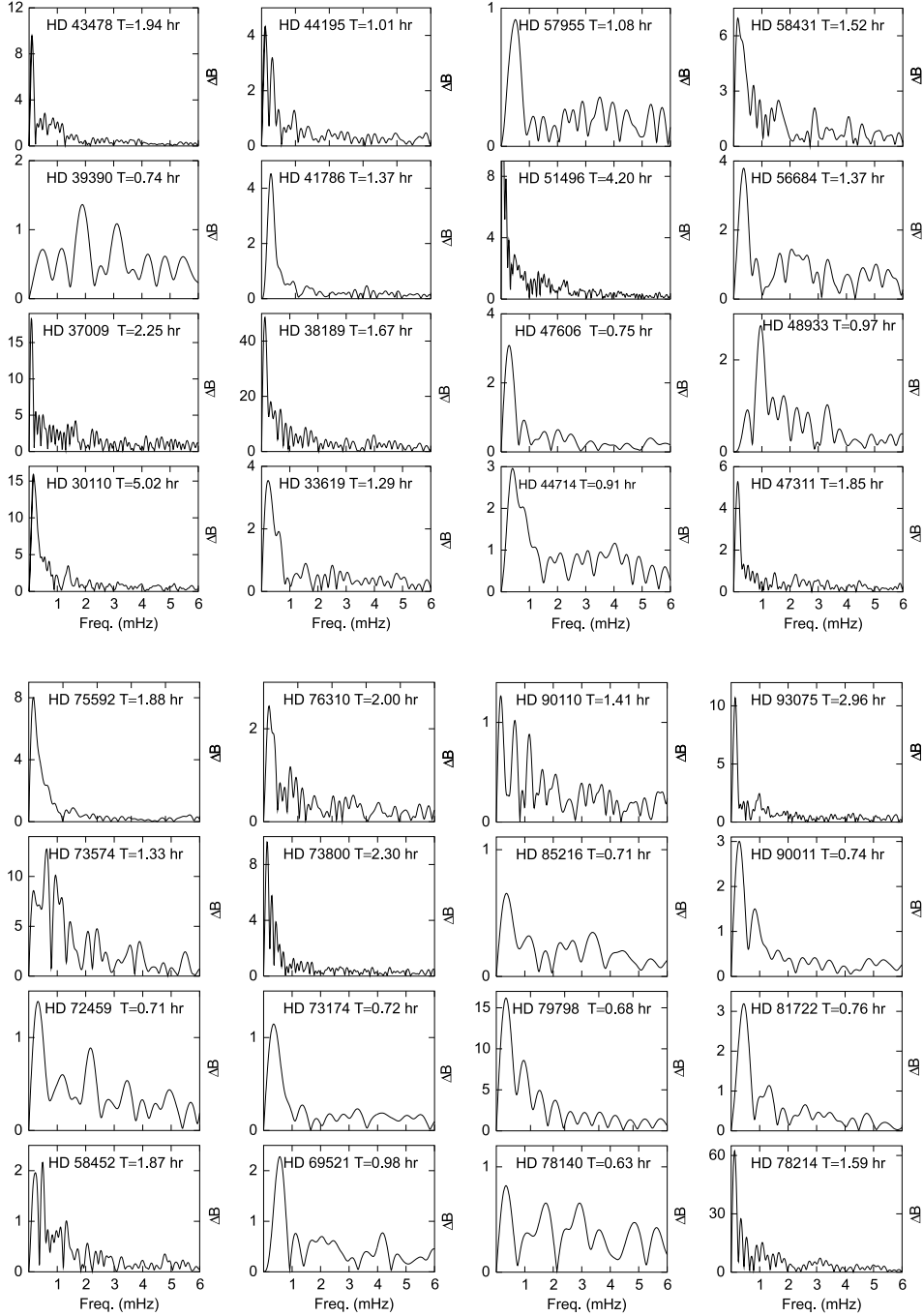


Figure 6. (Continued)

will have a great importance towards the field of asteroseismology of peculiar stars. HD 104202C is, therefore an interesting candidate for the ongoing Nainital–Cape survey for Ap and Am stars.

4. Null results

The amplitude spectra of the other 48 candidates are shown in Fig. 6. Each panel covers the frequency range 0–6 mHz and an amplitude range of several milli-magnitude, depending upon sky background and scintillation noise. We caution that the appearance of a star in our list of negative results does not exclude it from being a variable star. The star may have been non-variable for the duration of these particular observations, or the noise level was sufficiently high to hide the variability. Apart from this, it must be kept in mind that the peak corresponding to the main frequency in the roAp pulsation spectra can appear and disappear on a time scale of days (Handler 2004). Furthermore, the beating between close frequencies may interfere destructively to hide the variability. Hence, these results can be used as an informative reference, but all these stars should be observed again in future to check for variability. The Nainital–Cape survey is an on-going project. In addition to the main-sequence stars – Ap and Am stars – luminous peculiar stars have been included in our list of variable candidates, according to the “p” peculiarity criterion defined by Masana *et al.* (1998).

5. Conclusions

HD 98851 and HD 102480 are variables with unusually high overtones having nearly harmonic (or sub-harmonic) period ratios of 2:1. Two close pulsation periods in HD 13038 (28 min and 34 min), with a beat period of 2.78 hr have been observed. HD 13079 appears to be an Am star pulsating in the fundamental mode with a period of 78 min. A δ Scuti type variability was discovered in HD 11378, with a period of approximately 2.3 hr. All the negative results of our observations have been presented and can be used as an informative reference for further observations.

Acknowledgements

I am thankful to Ram Sagar, Peter Martinez and S. Seetha for their guidance and for providing constant encouragement and stimulating discussions throughout my thesis work. I thank D. W. Kurtz, S. K. Gupta, B. N. Ashoka and Girish for their friendly support. I thank Dr. D. Mary and Amitesh for reading this manuscript rigorously.

References

- Abt, H. A. 1984, *ApJ*, **285**, 247.
 Ashoka, B. N., Seetha, S., Raj, E., Chaubey, U. S., Gupta, S. K., Joshi, S., Martinez, P., Kurtz, D. W., Sagar, R., Kasturirangan, K. 2000, *BASI*, **28**, 251.
 Deeming, T. J. 1975, *Ap&SS*, **36**, 137.
 Elkin, V. G., Riley, J. D., Cunha, M. S., Kurtz, D. W., Mathys, G. 2005, *MNRAS* (in press).
 ESA 1997, The Hipparcos and Tycho Catalogue, European Space Agency, SP 1239, ESA Publications, Division, ESTEC, Noordwijk, The Netherlands.
 Handler, G. 2004, *Communication in Asteroseismology*, **145**, 71.
 Joshi, S., Girish, V., Martinez, P., Sagar, R., Ashoka, B. N., Gupta, S. K., Seetha, S., Kurtz, D. W., Chaubey, U. S. 2000, *IBVS*, 4900.
 Joshi, S., Girish, V., Sagar, R., Martinez, P., Seetha, S. 2002, *Communication in Asteroseismology*, **142**, 50.

- Joshi, S., Girish, V., Sagar, R., Kurtz, D. W., Martinez, P., Kumar, B., Seetha, S., Ashoka, B. N., Zhou, A. 2003, *MNRAS*, **344**, 431.
- Kurtz, D. W. 1985, *MNRAS*, **213**, 773.
- Martinez, P., Kurtz, D. W., Kauffmann, G. M. 1991, *MNRAS*, **250**, 666.
- Martinez, P., Ashoka, B. N., Kurtz, D. W., Gupta, S. K., Chaubey, U. S. 1999a, IBVS, 4677.
- Martinez, P., Kurtz, D. W., Ashoka, B. N., Chaubey, U. S., Gupta, S. K., Leone, F., Catanzaro, G., Sagar, R., Raj, E., Seetha, S., Kasturirangan, K. 1999b, *MNRAS*, **309**, 871.
- Masana, E., Jordi, C., Maitzen, H. M., Torra, J. 1998, *A&AS*, **128**, 265.
- North, P. 1993, In: *Peculiar Versus Normal Phenomena in A-type and Related stars*, IAU Colloq. No. 138, p. 577.
- Roman, N. G., Morgan, W. W., Eggen, O. J. 1948, *ApJ*, **107**, 107.
- Sagar, R. 1999, *Current Science*, **77**(5), 643.
- Seetha, S., Chaubey, U. S., Girish, V., Joshi, S., Ashoka, B. N., Gupta, S. K., Kurtz, D. W., Martinez, P., Sagar, R. 2001, *BASI*, **29**, 309.
- Smith, M. A. 1974, *ApJ*, **189**, 101.
- Stibbs, D. W. N. 1950, *MNRAS*, **110**, 395.
- Titus, J., Morgan, W. W. 1940, *ApJ*, **92**, 256.
- Zhou, A.-Y. 2001, *Communication in Asteroseismology*, **140**, 59.

HD 12098 and Other Results from Nainital–Cape Survey

V. Girish

Tata Institute of Fundamental Research, Homi Bhabha Road, Mumbai 400 005, India.

e-mail: giri@tifr.res.in

Abstract. Nainital–Cape Survey was started with an aim to search for new rapidly oscillating Ap stars in the northern hemisphere. We discovered one new mono-periodic roAp star HD 12098. The frequency separation of HD 12098 suggests a rotation period of 5.5 days for the star. We summarize here the observations of HD 12098 and briefly discuss the results of the multi-site observation campaign organized to resolve the ambiguity in the determination of the rotation period of HD 12098. Other interesting results like non-oscillating Ap stars discovered and two candidate stars in which roAp periodicity is seen but not confirmed are also discussed.

Key words. Stars: roAp stars—stars-individual: HD 12098—stars-individual: HD 207561—stars-individual: HD 17431—stars-individual: HD 25499—stars-individual: HD 38143—stars-individual: HD 38817.

1. Introduction

Nainital–Cape Survey is a program initiated with the main aim of detecting and studying new northern hemisphere rapidly oscillating Ap stars (roAp) stars. RoAp stars are short-period photometric variables with periods ranging from 5–21 minutes with typical amplitudes of few milli magnitude. Hence, the study of roAp stars requires a very stable site and atleast one meter class telescope. For these reasons, the 104 cm Sampurnanand telescope at ARIES, Nainital was selected for the survey. For a summary of the Nainital–Cape Survey, site characteristics and the facilities available at ARIES, Nainital for variable star research, see, Sagar & Mary (2005) in these proceedings.

Nainital–Cape Survey is an ongoing program. Till 2004, we have searched for short period variability in a total of 63 stars and detected rapid oscillations in HD 12098 (Girish *et al.* 2001) and δ Scuti oscillation in four stars (Joshi 2005). The selection of the candidate stars for the survey was made mainly on the basis of their Ap nature and Strömgren colours, and temperature whenever available. On the basis of known roAp stars, Matinez & Kurtz (1995) found that roAp stars fall in specific Strömgren colour space. To increase our chances of detecting new roAp stars, the candidates for the survey were selected mainly on these empirical limits on the six Strömgren colours. Since the limits are only empirical, the selection criteria is relaxed so as to include those stars which might lie just outside these limits. In fact one of the six Strömgren colours (δm_1) of HD 12098, the first roAp star discovered under the survey falls outside of the corresponding limit.

Ap stars which share similar properties as that of roAp stars, but do not show rapid variability are called as non-oscillating Ap (noAp) stars. The noAp stars hold equal importance as the roAp stars for the understanding of roAp phenomenon. During survey, we found that three stars HD 25499, HD 38143 and HD 38817 fall under the category of noAp stars. The amplitude of variability, if any, is less than 0.2 mmag in these three stars in the frequency range 1–5 mHz. In addition, two stars HD 17431 and HD 207561 showed variability on a few nights, but the data are not enough to confirm variability in these stars. These results are discussed in the following sections.

2. Rapidly oscillating Ap stars

2.1 HD 12098

HD 12098 is the first roAp star discovered under the Nainital–Cape Survey. The star is a bright ($m_V = 7.9$) F0 star. The Strömgren colour indices of the star ($b - y = 0.191$, $m_1 = 0.328$, $c_1 = 0.517$ and $\beta = 2.796$, Hauck & Mermilliod 1998) combined with the de-reddened parameters estimated using the calibrations by Crawford (1975) falls within the empirical limits suggested for roAp stars except for δm_1 . The discovery of roAp oscillations in HD 12098, hence extends the limits on δm_1 suggested by Martinez & Kurtz (1995) for probable roAp stars by a slight margin.

On the basis of Strömgren colours, HD 12098 was selected as a candidate for the survey and observed on the night of 21st November 1999. The discovery light curve plotted in Fig. 1 clearly shows variability around 7.1 minute. To confirm this variability, the star was observed on following nights.

The discovery and follow-up observations showed that the oscillation amplitude in HD 12098 varies from one night to the following night. This indicates plausible multi-mode oscillation and/or rotational modulation. To determine the nature of oscillation in HD 12098 it was observed on six nights continuously from Gurushikhar Observatory, Mt. Abu and Sampurnanand telescope, Nainital for a total of 65 hours.

The analysis of an individual nights' data clearly shows the modulation in the amplitude of oscillation. In roAp stars, the amplitude modulation can be caused either by the beating of close spaced periods and/or due to the rotation of the star explained by Oblique pulsator model (Kurtz 1982). To identify the nature of pulsations in HD 12098, we combined the data obtained from Mt. Abu on a common time scale and subjected

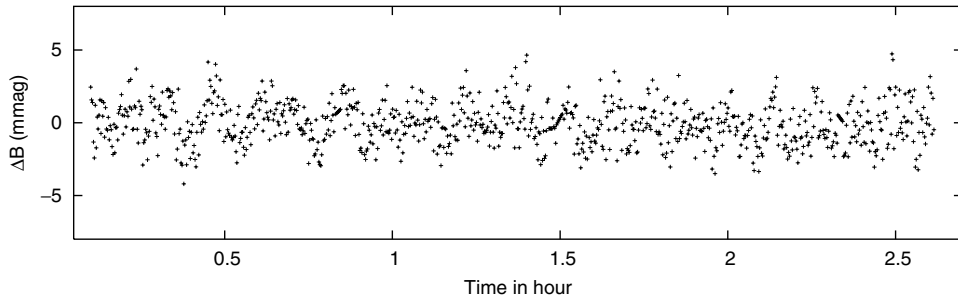


Figure 1. Discovery lightcurve of HD 12098 observed on the night of 21st November 1999. Each dot represents a 10-s integration.

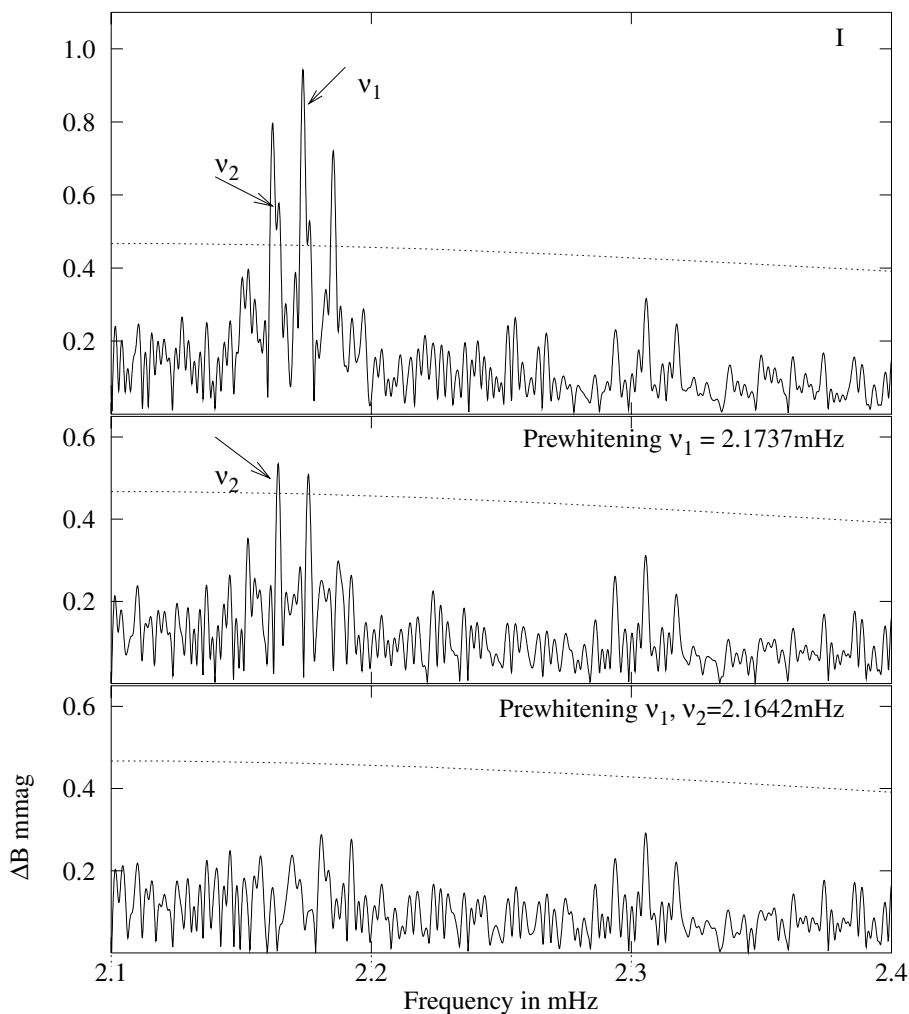


Figure 2. Amplitude spectra of the data obtained from Mt. Abu merged together. See text for details.

it to discrete Fourier transform (DFT). The resulting amplitude spectra shows a peak around 2.173 mHz. To avoid the identification of spurious peaks as real, the significance limit is set on the basis of local noise. The running average of amplitudes of a group of fifty frequencies are computed as the local noise assuming all the peaks in the data as noise. The significance level is set at four times the local noise level (see, Breger *et al.* 1996). Any peak lying above this limit is treated as being significant. In Fig. 2, this significance limit is shown as a dotted line.

In Fig. 2, we can see an unresolved frequency very close to the alias peak (marked ν_2). To recover this frequency and to check for any other peaks that might be buried under the main frequency at ν_1 , we subtracted a noise-free sinusoid with the amplitude and frequency corresponding to that of ν_1 from the time series. This process is typically referred to as prewhitening. The residual data are subjected to DFT again. The resulting

amplitude spectra is plotted in the second panel of Fig. 2. We can see two peaks at the 2.1641, 2.1759 mHz with almost the same magnitude. The frequency difference between the two corresponds to $1c/d$ suggesting that one of them is the alias of the second frequency. Typically the real frequency will be of a higher amplitude and the aliases will be of a lower amplitude. But, for low amplitude peaks near the noise level, this will not be true and it will not be possible to identify the real peak from the alias.

The second frequency ν_2 is interpreted as the rotationally split component of the main frequency ν_1 (Girish *et al.* 2001). Assuming that $\nu_2 = 2.1641$ mHz is the real frequency, the authors predict a rotation period of 1.22 day for HD 12098. Instead, if $\nu_2 = 2.1759$ mHz, then the rotation period will be 5.5 day. The magnetic field measurements (Wade *et al.* 2001) also favour the longer period.

2.2 Multi-site campaign

From the single site observations it is found that the second frequency is the rotationally split component of the oscillation frequency of HD 12098. However, the second frequency has a 1 cycle/day alias ambiguity.

With the main aim of resolving the ambiguity in determining the second frequency and hence, rotation period of the star, we organized a multi-site observation campaign on HD 12098 in October/November 2002. The campaign involved a total of eleven observatories with ten observatories finally contributing to the data. A total of 394 hours of useful data extending over 28 nights with 45% duty cycle were obtained.

From the analysis of the multi-site data, we obtained mainly five frequencies separated by $\delta\nu = 2.14 \pm 0.02$ mHz. We have already ruled out the plausibility of such a small difference arising due to independently excited modes (Girish *et al.* 2001). The equal separation of the frequencies seen in the multi-site data gives a rotation period of $\Omega = 5.41 \pm 0.05$ day for HD 12098. This is very close to the 5.5 day rotation period predicted by Girish *et al.* (2001) assuming $\nu_2 = 2.1759$ mHz as the real frequency instead of 2.1641 mHz. The amplitude modulation of the pulsation of the individual nights data also favours a period close to 5.41 day. The detailed observation and results of the multi-site campaign will be published elsewhere (Seetha S. *et al.* 2005).

3. Non-oscillating Ap stars

Non-oscillating Ap stars hold equal value in the understanding of the roAp phenomena. The systematic difference between the noAp stars and roAp stars will help us in identifying the reasons why some of the stars which share similar properties of roAp stars do not show oscillations. During our survey, we observed three Ap stars which fit in to noAp star category. We discuss these noAp stars briefly.

3.1 HD 25499

HD 25499 satisfies all the six Strömgren colours used for identifying probable roAp stars. The star was observed on six occasions for a total of nine hours and found to show no oscillations with an amplitude limit of 0.2 mmag in the frequency range of 1–5 mHz. The amplitude of oscillations if any, would be below 0.2 mmag at 3σ level. The lightcurve and the amplitude spectrum of HD 25499 observed on 10th November 2003 are plotted in Fig. 3 and Fig. 4 respectively.

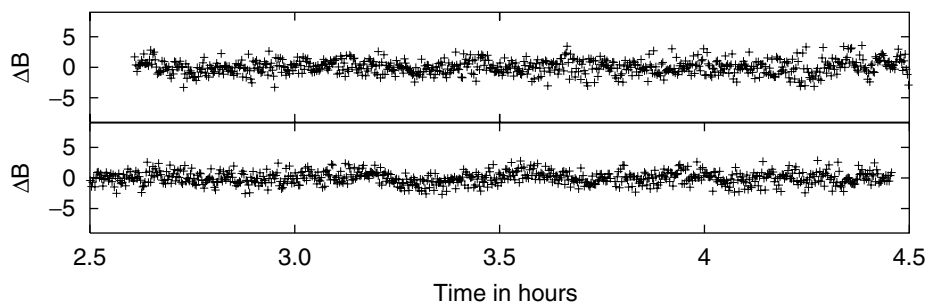


Figure 3. Lightcurve of HD 25499 observed on the night of 10th November 2003.

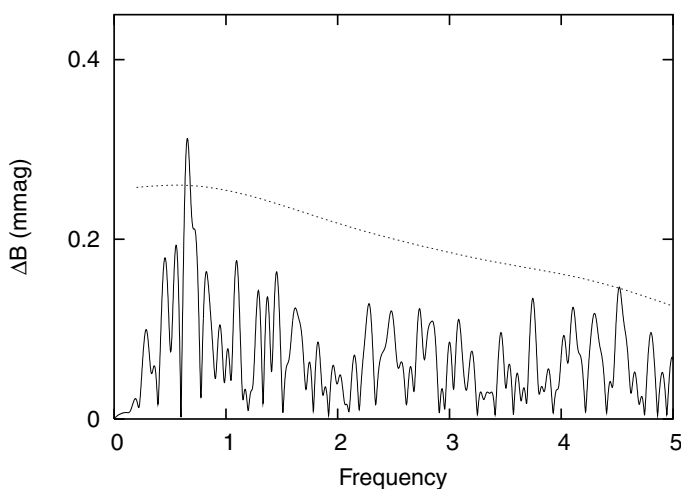


Figure 4. Fourier transform of the light curve of HD 25499 observed on 10th November 2003. The dashed curve represents three times the local noise level.

3.2 HD 38143

HD 38143 is classified as an A2 star whose Strömgren colour indices are well within the empirical limits suggested by Martinez and Kurtz (1995). The star was first observed on 23rd January 2000 and later, on three nights in 2002 and two nights in 2003 for a total of 22 hours. No variability is seen in all the runs with an amplitude limit of 0.3–0.4 mmag. The lightcurve of HD 38143 observed on 19th November 2003 is plotted in Fig. 5 with the corresponding amplitude spectra in Fig. 6. Our observations suggest that the star is a noAp star unless it has a very long rotation period of few years or the oscillations if any, have very small amplitudes (< 0.3 mmag).

3.3 HD 38817

HD 38817 is classified as an A2 star. Only three of the six Strömgren colours (H_β, m_1, dm_1) satisfy the limits suggested for a probable roAp star while one parameter, c_1 lies within 10% of the corresponding limiting value. The star was

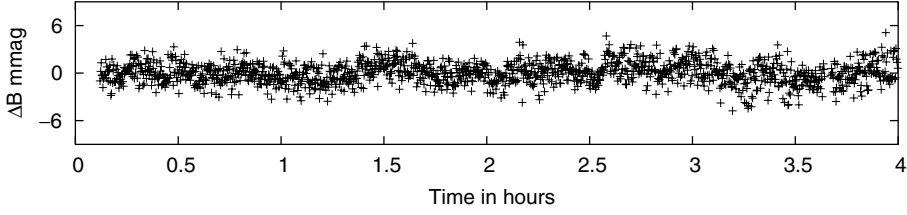


Figure 5. Amplitude spectra of HD 38143 observed on 19th November 2003.

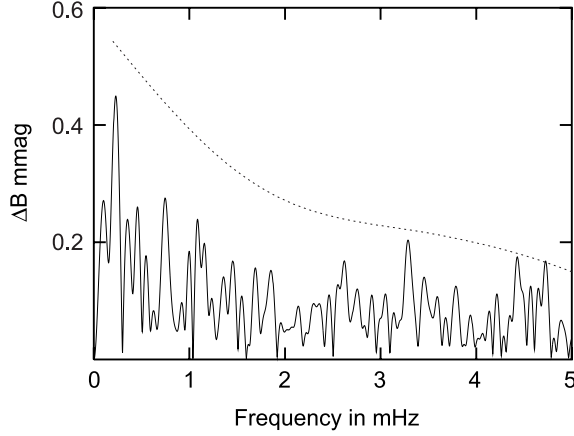


Figure 6. Amplitude spectra of lightcurve plotted in Fig. 5.

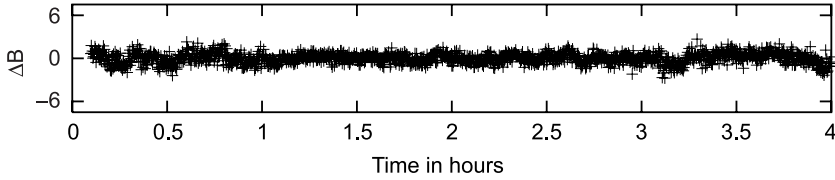


Figure 7. Lightcurve of HD 38817 observed on the night of 9th November 2003.

selected as a candidate star for Nainital–Cape Survey to check the possible extension of these limits, but turned out to be a non-variable at a limiting amplitude of 0.2 mmag. The star was observed for a total of 14 hours over four nights with the last three nights' observations extending for more than three hours each. The lightcurve and the corresponding amplitude spectrum of HD 38817 observed on 9th November 2003 are plotted in Figs. 7 and 8 respectively. From the analysis of the data we conclude that HD 38817 is a noAp star at a limiting amplitude of 0.4 mmag at 3σ confidence level.

4. Probable roAp stars

In the course of our survey, we observed variability in the frequency range of 1–5 mHz in two stars, HD 207561 and HD 17431. But, the data on hand are not sufficient to

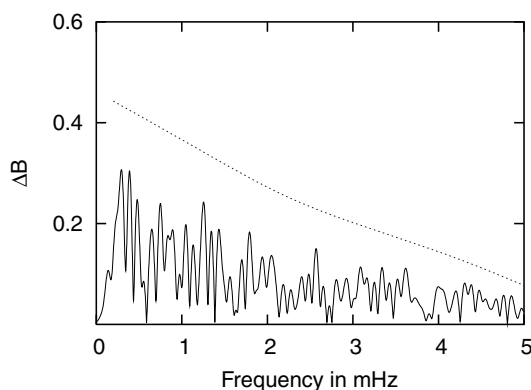


Figure 8. Plot of amplitude spectrum of the lightcurve plotted in Fig. 7.

confirm oscillations in these stars. HD 207561 is classified as an Am star and the detection of rapid variability will make it the first Am star showing roAp type variability. For the benefit of the others who may be interested in these stars, we briefly describe them.

4.1 HD 207561

HD 207561 is an F0 III star which falls well within the empirical limits for probable roAp stars (Martinez *et al.* 1995). The star was observed on 16th December 2000 for the first time and on several nights later. The lightcurve and the corresponding Fourier spectrum of HD 207561 obtained on two consecutive nights on 6th and 7th December 2002 plotted in Fig. 9 show regular variations which were absent in the comparison star data observed in the second channel. HD 207561 is classified as an Am star (Floquet 1975) and pulsations in this star if confirmed, will make HD 207561 the first Am star to show rapid oscillations similar to roAp stars.

Both the lightcurve and the amplitude spectrum of HD 207561 plotted in Fig. 9 clearly show regular variability around six minutes on two different nights. The absence of similar variation in the comparison star rules out the possibility of local effects. But the absence of the six minute period in the follow-up observations stops us in confirming rapid oscillations in the star. Also the period is very close to six minutes. Whenever periodicity of the integral multiple of a minute is detected, care should be taken to rule out the plausibility of drive error or other instrumental effects mimicking regular variability. Another likely reason for the non-detection of oscillations in the follow-up nights might be due to mis-identification of the star. However, the confirmation of the presence or absence of oscillations in HD 207561 will be important as it will make HD 207561 the first Am star with roAp type variability.

4.2 HD 17431

HD 17431 is classified as an A3 star. The Strömgren colour indices of the star satisfy all the six Strömgren parameters suggested for a probable roAp star. The star was first observed on the night of 9th October 2000. The Fourier transform of the data showed a

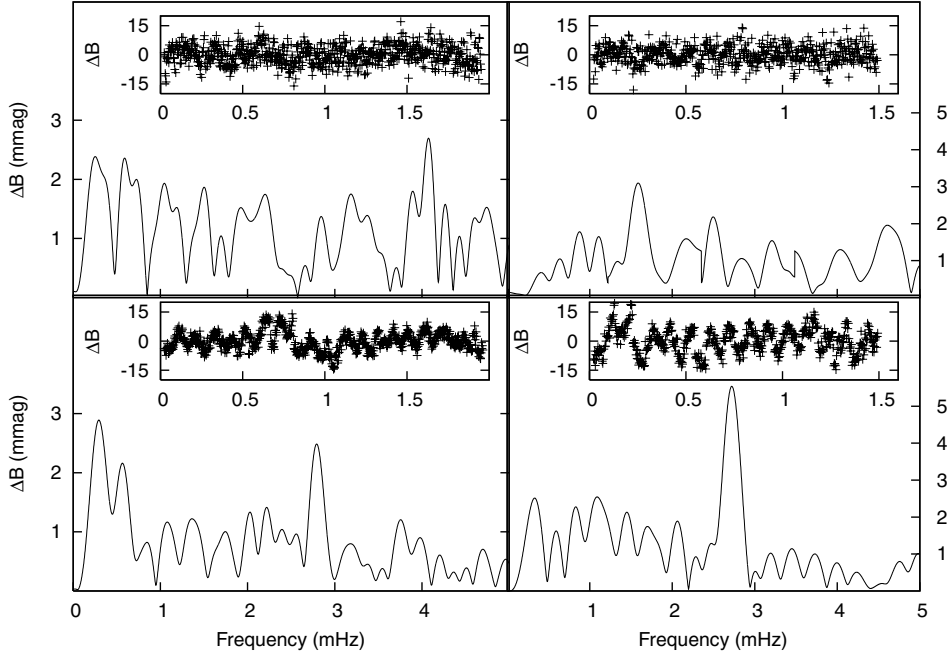


Figure 9. Amplitude spectrum of HD 207561 observed on the nights of 6th December 2000 (left panels) and 7th December 2000 (right panels). The top two panels show the plots of the comparison stars while the bottom panels correspond to HD 207561 data. A peak around 6.1 minute period is clearly seen in both amplitude spectrum and lightcurve of HD 207561 which is absent in the comparison star.

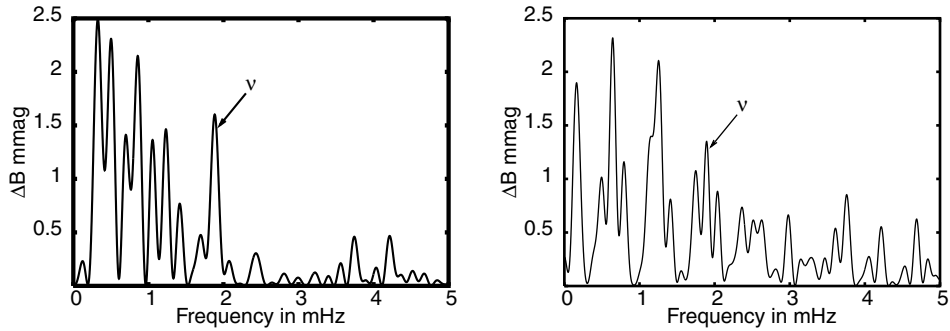


Figure 10. Amplitude spectrum of HD 17431 observed on 10th October 2000 and 6th December 2001. The peaks marked as ν correspond to ~ 8.8 minutes.

peak around 8.8 minutes. A similar period was observed on the follow-up observations though not with very good signal except on 10th October 2000. In Fig. 10, we plot the amplitude spectrum of the data obtained on 10th October 2000 and 6th December 2001 which indicate the presence of a period around 8.8 minutes. Though, similar variability near this period is seen on few other nights, the nights were not photometric which prevents us from making any statement on this star.

5. Conclusions

We presented a brief summary of the main results of the Nainital–Cape Survey conducted from late 1999 to early 2004. During this period we have discovered one roAp star HD 12098 which is a single periodic oscillator. The amplitude modulation of the pulsation frequency and frequency splitting observed in HD 12098 suggest a rotation period of ~ 5.4 days for the star. We classify three stars HD 25499, HD 38143 and HD 38817 as non-oscillating Ap stars which can be useful in understanding the differences between roAp and Ap stars which share similar properties of roAp stars but do not oscillate. We also presented two interesting objects HD 17431 and HD 207561 which showed variability on few occasions. Further observations on the two may confirm the roAp variability in the two. The confirmation of oscillations in HD 207561 may be a very useful result as it will be the first Am star showing roAp type variability.

Acknowledgements

The author would like to acknowledge Dr. S. Seetha for the opportunity to work in the exciting field of roAp stars and who was also instrumental in organizing the multi-site campaign on HD 12098. Thanks are due to Prof. Ram Sagar for all the support at ARIES, Nainital. The author would like to acknowledge the help and support of Dr. Santosh and all the other observatory staff of ARIES, Nainital without whose help this survey would have been impossible.

References

- Breger, M., Handler, G. and other 20 co-authors 1996, *A&A*, **309**, 197.
Crawford, D. L. 1975, *AJ*, **80**, 955.
Floquet, M. 1975, *A&AS*, **21**, 25.
Girish, V., Seetha, S., Martinez, P., Joshi, S., Ashoka, B. N., Kurtz, D. W., Chaubey, U. S., Gupta, S. K., Sagar, R. 2001, *A&A*, **380**, 142.
Hauck, B., Mermilliod, M. 1998, *A&AS*, **129**, 431.
Joshi, S. 2005, this volume.
Kurtz, D. W. 1982, *MNRAS*, **200**, 807.
Martinez, P., Kurtz, D. W. 1995, *AP&SS*, **230**, 29.
Ram Sagar, Mary, D. L. 2005, this volume.
Seetha, S., Girish, V. *et al.* 2005 (in preparation).
Wade, G. A., Bagnulo, S., Donati, J. F., Lueftinger, T., Petit, P., Sigut, T. A. A. 2001, *ApN Nu.* 35.

Asteroseismic Theory of Rapidly Oscillating Ap Stars

Margarida S. Cunha

Centro de Astrofísica da Universidade do Porto, Rua das estrelas, 4150, Porto, Portugal.

e-mail: mcunha@astro.up.pt

Abstract. This paper reviews some of the important advances made over the last decade concerning theory of roAp stars.

Key words. Stars: variable, chemically peculiar, roAp, magnetic.

1. Introduction

The success of the asteroseismic studies of rapidly oscillating Ap (roAp) stars depends strongly on our ability to understand their oscillation spectra. Questions like: *which modes are excited and why, what is the expected spacing between eigenfrequencies, how many components are expected to be found in the multiplet structures, or what are the relative amplitudes of the different components of the multiplets* have been central to recent theoretical studies of roAp stars. In this paper, we will review different aspects of the theoretical work recently carried out in this field.

2. From spherical symmetry to roAp stars

A key aspect to keep in mind when studying pulsations in roAp stars is the presence of different physical agents which influence the pulsations in a non-spherically symmetric way. The deviations from spherical symmetry add additional complexity to the interpretation of the oscillation spectra of these stars, as well as to their asteroseismic study in general. Through the rest of this section we will attempt to summarize some of the properties of the eigenfrequencies and eigenfunctions of modes of oscillation with fixed radial order n and fixed degree l , in models with different symmetry properties, without going through the cumbersome calculations needed to derive them. More information on the subject can be found in Unno *et al.* (1989) and Gough (1993), and, for the specific case of roAp stars, in Bigot & Dziembowski (2002) and Gough (2003). We will consider linear, adiabatic pulsations. Moreover, we will assume that the impact of the symmetry-breaking agents on pulsations is sufficiently small that their effects can be superposed linearly. Some of the non-axisymmetric agents present in roAp stars, particularly the magnetic field, can distort the modes of oscillation to an extent that a single eigenmode will no longer be associated to a single value of l (Dziembowski & Goode 1996). However, the implications of this fact will be discussed in the next section. Thus, for the rest of this section it will be assumed that each eigenmode is well represented by a single degree value l .

The modes of oscillation in the models that will be considered are described by the eigenvalue equation $\omega^2 \psi = \mathcal{A}(\psi)$ where ω are the eigenfrequencies, ψ are scalar

eigenfunctions (e.g., radial component of the displacement) and $\mathcal{A}$ is a generic differential operator which takes a different form depending on the model considered.

In a spherically symmetric model the choice of direction for the polar axis of our coordinate system is arbitrary: all equilibrium quantities in the differential operator $\mathcal{A}$ are independent of that choice. Since the orientation does not influence the eigenvalue problem, the $2l + 1$ eigenfrequencies associated with a mode of radial order n and degree l will be degenerate. The angular parts of the corresponding $2l + 1$ eigenfunctions are described by arbitrary (independent) linear combinations of the $2l + 1$ spherical harmonics Y_l^m (with the azimuthal order m varying from $-l$ to l). In particular, one can define a set of $2l + 1$ independent eigenfunctions ψ_m^0 for the spherical symmetric model, such that each is proportional to a different spherical harmonic function Y_l^m , with l fixed.

Next consider the case in which the spherical symmetry of the problem is broken due to an axisymmetric effect produced by a given physical agent, such as an axisymmetric magnetic field, or rotation. In this case $\mathcal{A}$ incorporates the spherically symmetric differential operator referred to above plus additional terms associated with the direct and indirect effects of the non-spherically symmetric agent on the oscillations. Because of the lack of spherical symmetry, the orientation of the eigenmodes is no longer arbitrary. If we use the axis of symmetry of the symmetry-breaking agent as the polar axis ($\theta = 0$), in a spherical coordinate system (r, θ, ϕ) attached to the star, the additional effects incorporated in $\mathcal{A}$ will be independent of ϕ . We can then have two situations: if the non-axisymmetric effect does not depend on the sense of the axis of symmetry (e.g., the Lorentz forces generated by the perturbed magnetic field), then the problem is invariant under reflection about the equator and some degeneracy remains. If, on the other hand, the effect depends on the sense of the axis of symmetry (e.g., the Coriolis force produced by the rotation of the star), then the degeneracy is completely lifted. In the first case, all but one of the $2l + 1$ eigenfrequencies will be degenerated in pairs, and the angular part of the corresponding pairs of eigenfunctions will be arbitrary (independent) linear combinations of the pairs of functions Y_l^{-m} and Y_l^m . Moreover, the angular part of the eigenfunction associated with the ‘unpaired’ eigenfrequency will be given by Y_l^0 . In the second case, the $2l + 1$ eigenfrequencies will all be different. Moreover, the arbitrariness in the angular part of the eigenfunctions will be totally removed: the $2l + 1$ eigenfunctions will each have an angular part described by a spherical harmonic function Y_l^m of a different azimuthal order m . Thus, if the perturbation is axisymmetric, we can still define a set of $2l + 1$ independent eigenfunctions ψ_m^1 that are solutions to the problem, each proportional to a different spherical harmonic function Y_l^m .

In roAp stars there are several symmetry-breaking agents which simultaneously influence the oscillations. This is the case of rotation, magnetic fields and possible structural differences associated with the spots observed at the surface of these stars. The fact that the axes of symmetry of these agents (assuming they are all axisymmetric), are not generally aligned, increases the complexity of the problem. To proceed with this description, let us specify the symmetry-breaking agents in our problem, by adopting the traditional oblique rotator model: a model of a magnetic rotating star, in which the axis of symmetry of the axisymmetric magnetic field is inclined in relation to the rotation axis.

If the magnetic field is force free (as is commonly assumed in the theoretical works of roAp stars), it will essentially affect the oscillations directly, via the Lorentz force,

associated with the perturbed magnetic field. Moreover, according to Bigot & Dziembowski (2002) the most important effects of rotation regarding the pulsations of roAp stars are the effect of the Coriolis force and the effect of the centrifugal distortion of the star. In fact they point out that, due to the high radial order of the oscillations observed in roAp stars, the latter, despite being of second order in the angular velocity Ω , is generally larger than the former. An essential difference between the effect of the Coriolis force and the effects of the Lorentz force and centrifugal distortion, is that only the first of these can detect the sense of the corresponding axis of symmetry. As shown in Bigot & Dziembowski (2002), this property of the Coriolis force has important observational consequences, even when the effect of the latter is small compared to the other two effects.

Since the axes of symmetry of the effects considered are not aligned, if we chose our coordinate system such that the polar axis is aligned with one of these axes, the effect produced by the other symmetry-breaking agent will depend on both θ and ϕ . Hence, the angular parts of the $2l + 1$ eigenfunctions will be given by different (but not arbitrary) linear combinations of the $2l + 1$ functions Y_l^m defined in that coordinate system. Generally there will be no degeneracy in the solutions. The $2l + 1$ eigenfrequencies will be all different and the linear combinations that define the angular parts of the eigenfunctions will be all well determined. Hence, unlike before, it is generally not possible to associate each eigenfunction with an individual value of m in either of the two coordinate systems.

To simplify our discussion further, let us consider the specific case of dipole modes (*i.e.*, $l = 1$). If we were to neglect the effect of the Coriolis force, then the problem would be invariant under the transformation $m \rightarrow -m$. Thus, the coefficients of the terms Y_1^1 and Y_1^{-1} in the linear combinations that define the angular part of the eigenfunctions would necessarily have the same absolute value. In fact, if there were no Coriolis force, a coordinate system would exist (with polar axis indicated by P , in Fig. 1) such that one of the eigenfunctions would be axisymmetric about the polar axis (*i.e.*, its angular part would be proportional to Y_1^0) while the other two eigenfunctions would each be axisymmetric about one of the other two axis of the coordinate system

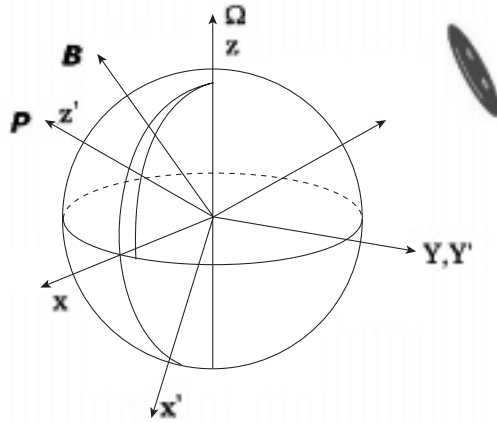


Figure 1. Schematic view of the coordinate systems that are most relevant to the study of pulsations in roAp stars (see text).

(*i.e.*, their angular parts in that coordinate system would be proportional, respectively, to $Y_1^{-1} + Y_1^1$ and $i(Y_1^{-1} - Y_1^1)$). The polar axis of that coordinate system is in the plane defined by the magnetic and rotation axes, and its position in that plane depends on the relative strengths of the magnetic and centrifugal effects.

The effect of the Coriolis force on the pulsations of roAp stars is typically small compared with the magnetic and the centrifugal effects. However, unlike the other effects, the former generates an asymmetry between the absolute values of the $m = 1$ and $m = -1$ coefficients of the linear combinations that determine the eigenfunctions (Bigot & Dziembowski 2002). As a consequence, the pulsation axes of the three eigenfunctions, which in the absence of the Coriolis force would be directed along the axes of the reference frame described above, change direction, over the period of pulsation, in planes which are perpendicular to the plane defined by the rotation and magnetic axes. The maximum of the radial displacement vector of each eigenmode describes an ellipse, over the corresponding period of pulsation, in the corresponding plane.

When the Coriolis effect is not neglected the eigenmodes can alternatively be described in a precessing reference frame (Gough 2003, 2005). The polar axis of that reference frame is slightly deviated from the plane defined by the rotation and magnetic axes and precesses about the polar axis of the pulsation coordinate system defined in the absence of the Coriolis effect (P in Fig. 1), on a long time scale. The eigenmodes of the problem defined in this precessing frame correspond to specific combinations of the eigenmodes that are solutions to the problem in the reference frame attached to the star. In the precessing reference frame one of the eigenmodes will be axisymmetric about the polar axis while the other two linearly independent modes will precess in a plane perpendicular to it.

3. The observer's view

Viewed from the earth the eigenfunctions will be proportional to linear combinations of the type $\sum_{m=-l}^l A_{l,m} \cos(\omega - m\Omega)t$, where Ω is the angular velocity of the star and t is the time. Thus, each of the $2l + 1$ eigenmodes will generally be seen as a $(2l + 1)$ -component multiplet. Because of the effect of the Coriolis force, the coefficients $A_{l,m}$ and $A_{l,-m}$ will generally have different absolute values. This inequality of the side peaks associated with the same value of $|m|$ depends on the relative importance of the Coriolis force and on the orientation of the mode plane. Moreover, the ratio between the sum of the pairs of side peaks to the central peak depends both on the inclination of the rotation axis to the line of sight and on the orientation of the mode plane.

As mentioned in the previous section, the magnetic field can distort the modes of oscillation to an extent that a single mode might no longer be well represented by a single spherical harmonic of degree l . Even though the overall effect of the magnetic field on the oscillations is small, there is a region, near the surface of the star, where the magnetic effect cannot be regarded as a small perturbation. It is this region, where the magnetic pressure is comparable or larger than the gas pressure, that is responsible for the significant distortion of the eigenmodes. In the previous section, this distortion of the eigenfunctions from single spherical harmonics was neglected. However, the latter has been studied by several authors (*e.g.*, Dziembowski & Goode 1996; Saio & Gautschi 2004) in a context in which only the magnetic effect was accounted for (*i.e.*, the effect of rotation on the oscillations was neglected). In this case, the angular part of the eigenfunctions is represented, in a coordinate system with

the polar axis along the magnetic axis, by sums of spherical harmonics of different degree l and the same azimuthal order m , *i.e.*, $\sum_l B_l Y_l^m$. In the observer's reference frame, each of the individual l components of the linear combination that describes the angular part of a mode of frequency ω will give rise to a $(2l + 1)$ -component multiplet of the type $\sum_{m=-l}^l b_{l,m} \cos(\omega - m\Omega)t$. Thus, as seen by the observer, an eigenmode will be associated with a $(2l_{\max} + 1)$ -component multiplet, where $l_{\max}$ is the maximum l value such that the corresponding coefficient $B_{l_{\max}}$, taken at the surface of the star, is still sufficiently large for that component to be observable after being averaged over the visible disk. Note that because the rotational effect on the oscillations has been neglected in these studies, the absolute value of the coefficients $b_{l,m}$ and $b_{l,-m}$ obtained in the latter is the same.

4. Basic signatures

In order to interpret the power spectra of roAp stars we need also to understand how the eigenfrequencies are modified by the non-axisymmetric effects described in the previous sections. In particular, it is most relevant to learn how basic signatures used in asteroseismic studies, like the small and large separations are affected. In fact, the large separations of roAp stars have been compared to theoretical predictions by Matthews *et al.* (1999) and Cunha *et al.* (2003), using standard spherically symmetric models, and in both cases discrepancies between theory and observations were found. Cunha *et al.* (2003) have attempted to modify some of the parameters included in the models used for a particular roAp star (like chemical abundance and amount of overshooting in the core), but the best agreement the authors found implied values for the metal abundance that were too low to be acceptable. Even though they could reconcile the theoretical results with the observations when the error-bars in the luminosity and effective temperature of the star were taken into account, the fact that the discrepancy between the theoretical predictions and the observations is systematic for roAp stars (Matthews *et al.* 1999) might indicate that some important physics was missing in the spherically symmetric models used.

The effect of the magnetic field on the frequency of the oscillations in roAp stars has been studied by different authors who accounted for the fact, first noted by Biront *et al.* 1982), that in the surface layers, where the gas pressure is low, the magnetic field effect cannot be treated as a small perturbation (e.g., Dziembowski & Goode 1996; Bigot *et al.* 2000; Cunha & Gough 2000; Cunha 2001; Saio & Gautschi 2004).

In the region where the gas and magnetic pressures are comparable, the waves are magnetoacoustic. Where the gas pressure dominates, however, the latter decouple into magnetic and acoustic components. Due to the rapid increase of the magnetic wavenumber, as the density gets larger, it is expected that the magnetic waves will dissipate (Roberts & Soward 1983). Therefore, the acoustic wave in the interior is continuously losing energy through the coupling that takes place in the magnetic boundary layer (see Fig. 2). As a consequence of this energy loss, the eigenfrequencies are complex, even when the oscillations are studied in an adiabatic approximation and are fully reflected at the surface. Figure 3 shows the imaginary part of the frequency perturbations, which originate from the coupling described above, in a polytropic model of an roAp star. It is clear from the figure that there are frequency intervals for which the dissipation is very large, when compared with other frequency intervals for which the latter is negligible. This fact might influence the frequency range in which

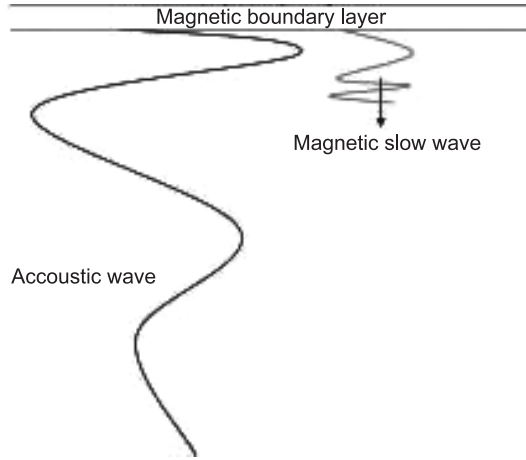


Figure 2. Schematic representation of the magnetic effect on the oscillations of roAp stars. There is a region in the magnetic boundary layer where the waves are magnetoacoustic. In the interior, the magnetic and acoustic parts decouple and the magnetic component is expected to dissipate due to the rapid increase in its wavenumber.

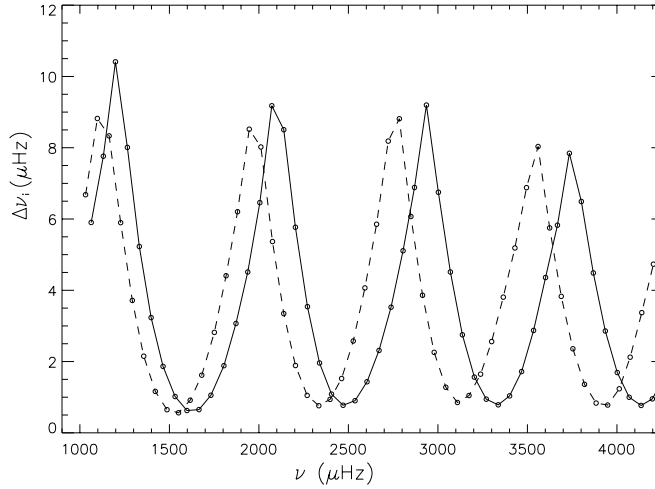


Figure 3. Imaginary part of the magnetic perturbations to the oscillations in two polytropic models of an roAp star: solid line – mass $M = 2.0M_{\text{sun}}$ and radius $R = 1.9R_{\text{sun}}$; dashed line – mass $M = 1.5M_{\text{sun}}$ and radius $R = 1.75R_{\text{sun}}$.

oscillations are excited in a given roAp star. Thus, magnetic coupling might introduce mode selection in some roAp stars.

Due to the high radial order of the oscillations observed in roAp stars, one might expect, to first approximation, to find the eigenfrequencies regularly spaced in the oscillation spectra. The large separations are defined as the difference in frequency between modes of the same degree l and consecutive radial orders, n . Thus, if modes of both even and odd degree are excited, the large separations will correspond to approximately twice the average spacing between consecutive modes in the oscillation

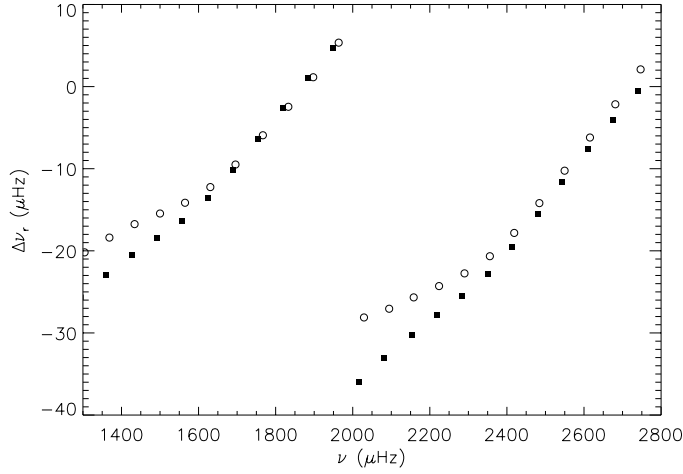


Figure 4. Real part of the magnetic perturbations to the oscillations of a polytropic model with a mass $M = 1.5M_{\text{sun}}$ and a radius $R = 1.75R_{\text{sun}}$. The open circles and filled squares show the perturbations for modes of degree $l = 1$ and $l = 3$, respectively.

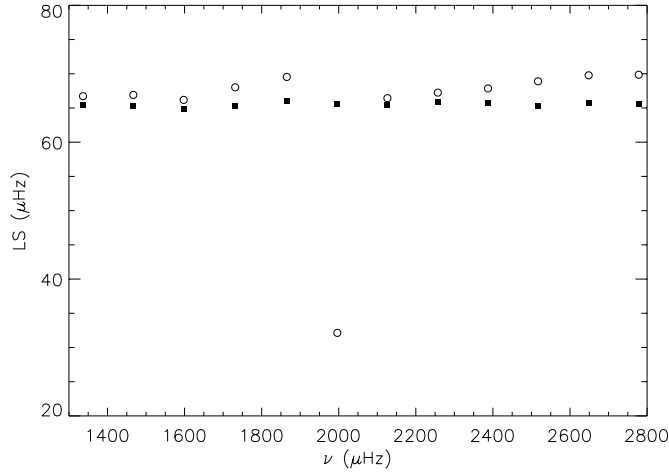


Figure 5. Magnetic perturbations to the large separations of the model described in the caption of Fig. 4. The perturbed large separations (shown by open circles) are compared with the unperturbed ones (shown by filled squares).

spectra. Moreover, modes of degree l , and $l + 2$ and, respectively, radial orders n and $n - 1$ are almost degenerated in frequency. The difference between their frequencies is known as the small separations and in roAp stars they are typically only a few μHz .

In Fig. 4, the real part of the magnetic perturbations to the eigenfrequencies of a polytropic model of a roAp star are shown for modes of degree $l = 1$ and $l = 3$. The dependence of the effect on the degree of the mode is clearly seen in some frequency domains. Thus the small separations will be significantly modified from the value they would have in the absence of a magnetic field. In Fig. 5, the perturbations

to the eigenfrequencies of the same model are combined, for the mode of degree $l = 1$, to show the magnetic perturbations to the large separations. As can be seen, the large separations are generally slightly increased by the effect of the magnetic field. However, it is also apparent from the figure that at some particular frequencies the large separations can be significantly modified. Thus, when inspecting the oscillation spectra of some roAp stars one might expect to find anomalies in the usual mode spacing, resulting from these abrupt changes in large separations at given frequencies.

5. Conclusion

The theory of rapidly oscillating Ap stars has progressed considerably over the past decade. Some of the major advances concern the effect of the magnetic field on the oscillations, and the combined magnetic and rotational effects on the orientation of the modes. As argued in this review, both of these issues have important implications to the interpretation of the oscillation spectra of roAp stars. However, further progress is desirable, particularly in what concerns a consistent theory that takes into account the combined effect of the magnetic field and rotation, without neglecting the distortion of the eigenmodes from single spherical harmonics. Additionally, progress is needed in what concerns the modelling of the mechanism by which the oscillations are reflected near the surface of roAp stars. Approximations like introducing a mechanical boundary condition that assures full reflection of the oscillations near the surface, or assuming a locally uniform magnetic field, can disguise the true reflection process. Since the large separations depend on the size of the pulsation cavity, further improvements on the modelling of the layers where reflection takes place would also be important.

This paper was centred on adiabatic studies of pulsations in roAp stars. However, progress has also been made concerning important theoretical issues that relate to non-adiabatic studies. In particular, it is believed today that oscillations in roAp stars are excited by the κ -mechanism in the hydrogen ionization region (Dziembowski & Goode 1996). High frequency oscillations are found unstable in models of roAp stars if the surface convection is assumed to be suppressed by the magnetic field at least in some angular region of the star (Balmforth *et al.* 2001; Cunha 2002), or if it is assumed that the star has a chromosphere (Gauthschiy *et al.* 1998). A theoretical instability strip is now available (Cunha 2002), which can be used to compare with the observations, as well as to motivate further observations of Ap stars that populate regions of luminosity and effective temperature where up-to-date no high frequency oscillations were found.

Acknowledgements

MC is supported by FCT-Portugal, through the grants BPD/8338/2002 and POCTI/CTE-AST/57610/2004.

References

- Balmforth, N. J., Cunha, M. S., Dolez, N., Gough, D. O., Vauclair, S. 2001, *MNRAS*, **323**, 362.
- Biront, D., Goossens, M., Cousens, A., Mestel, L. 1982, *MNRAS*, **201**, 619.
- Bigot, L., Dziembowski, W. A. 2002, *A&A*, **391**, 235.
- Bigot, L., Provost, J., Berthomieu, G., Dziembowski, W. A., Goode, P. R. 2000, *A&A* **356**, 218.

- Cunha, M. S. 2001, *MNRAS*, **325**, 373.
Cunha, M. S. 2002, *MNRAS*, **333**, 47.
Cunha, M. S., Fernandes, J. M. M. B., Monteiro, M. J. P. F. G. 2003, *MNRAS*, **343**, 831.
Cunha, M. S., Gough, D. O. 2000, *MNRAS*, **319**, 1020.
Dziembowski, W. A., Goode, P. R. 1996, *ApJ*, **458**, 338.
Gautschy, A., Saio, H., Harzenmoser 1998, *MNRAS*, **301**, 31.
Gough, D. O. 1993, *Astrophysical fluid dynamics*, Elsevier, Amsterdam, p. 399.
Gough, D. O. 2003, *ASP Conf. Series*, **305**, 389.
Gough, D. O. 2005, *The Roger Tayler Memorial Lectures*, Special issue of *Astronomy & Geophysics*, Royal Astronomical Society, p. 16.
Matthews, J. M., Kurtz, D. W., Martinez, P. 1999, *ApJ*, **511**, 422.
Roberts, P. H., Soward, A. 1983, *MNRAS*, **205**, 1171.
Saio, H., Gautschy, A. 2004, *MNRAS*, **350**, 485.
Unno, W., Osaki, Y., Ando, H., Saio, H., Shibahashi, H. 1989, *Nonradial Oscillations of Stars*, University of Tokyo Press, Japan.

The Dushak–Erekdag Survey of roAp Stars

Tatyana Dorokhova¹ & Nikolay Dorokhov

Astronomical Observatory, Odessa National University, Ukraine.

¹*e-mail: tnd@te.net.ua*

Abstract. The search of roAp stars at Mt. Dushak–Erekdag Observatory was started in 1992 using the 0.8 m Odessa telescope equipped with a two-star high-speed photometer. We have observed more than a dozen stars so far and discovered HD 99563 as roAp star while BD + 8087 is suspected to have rapid oscillations. Negative results of our observations for the search of rapid oscillations in four stars in NGC 752 are also discussed.

Key words. Stars: chemically peculiar—stars: oscillations—stars: variables: other—techniques: photometric.

1. Introduction

The first rapidly oscillating Ap (roAp) star was discovered by Kurtz in 1978 at the South-African Astronomical Observatory (SAAO). Details of the discovery and further studies are given in Kurtz (1990) and Kurtz & Martinez (2000). Some roAp stars have multiple modes of pulsations and they are valuable for asteroseismological studies.

Out of the 32 known roAp stars so far, about 84% were discovered in the southern hemisphere from SAAO by Kurtz (1990), Martinez *et al.* (1991) and Martinez & Kurtz (1994a, b). In order to investigate the cause for this asymmetric distribution in the two hemispheres of the sky, a number of surveys were carried out for discovering northern roAp stars (e.g., Matthews & Wehlau 1985; Heller & Kramer 1988; Nelson & Kreidl 1993). Until 1998 only two northern roAp stars were discovered. Later, the Vienna Survey (Handler & Paunzen 1999) and the Nainital–Cape Survey (Girish *et al.* 2001) discovered two more northern roAp stars, one by each group.

We commenced our survey programme for roAp stars in 1992 using the Odessa 0.8 m telescope with a two-star high-speed photometer (Dorokhov & Dorokhova 1994a) at the Mt. Dushak–Erekdag Observatory, Copet Dag Mountains, Central Asia, Turkmenistan. We started with the test observations of the well-studied roAp star HR 1217 (Dorokhov *et al.* 1994b). The results showed that the seeing at the observatory site and the instrumentation quality meet the precision requirements for photometry of roAp stars. The next section describes criteria used for the candidate selection. Details of the observations and the results obtained so far are given in the remaining part of the manuscript.

2. Candidate selection

Martinez *et al.* (1991) presented the detailed searching scheme based on Strömgren photometry. We used initially their photometric criteria and extracted the list of

Table 1. List of candidate Ap stars observed to detect rapid oscillations.

Star	m_v	Field region		Sp	Sp (Rn)
		α (2000)	δ (2000)		
HD 15257	5.38	2 28.2	+29 40	F0	
HD 17317	8.3	2 47.5	+21 21	A5p	
HD 99563	8.17	11 27.3	−08 52	F0	F0 Sr
HD 115606	8.3	13 18.0	+13 00	A2	A2 Sr
HD 217401	8.0	23 00.3	+14 00	A2p	A2 Sr
HD 276625	10.0	04 35.5	+41 08	A7	A7 CrEu
HD 281367	9.95	03 59.2	+32 43	A7	A8 SrEu
HD 281886	8.9	04 17.6	+31 26	F0p	F0 Sr
BD +8087	9.64	02 49.4	+81 26	Ap	F0 SrEu
NGC 752 cluster region					
R142, BD+36363	10.6			F2V	
R145, BD+36365	10.4			F0III	
R160, BD+37431	10.5			F2III	
R201	11.1			F3	

candidates from $uvby\beta$ Catalogue (Hauck & Mermilliod 1990, 1998). Later on, we also used the catalogues of Renson (1991); Lebedev (1986); Klochkova & Kopylov (1986), etc. for identifying Sr, Cr, Eu peculiarities in the stars. Candidates were selected from both field and cluster regions. Table 1 gives a list of stars which were chosen as candidates for observation. This table lists HD numbers, visual magnitudes, positions, spectral types from HD catalogue and spectral types from catalogue by Renson (1991) for all the observed stars. Star numbers provided by Rebeiro (1970) in NGC 752 are given along with BD catalogue number. Searching for roAp stars in clusters is an interesting idea. Asteroseismology of cluster stars presents fundamental tests to the theories of stellar structure and evolution. Similar attempts have been undertaken earlier (e.g., Matthews *et al.* 1988) but without success hitherto. NGC 752 provides an ideal test of a variety of evolutionary phenomena and it seems suitable due to the fact that it contains plenty of early F type stars. This intermediate-age (~ 1.8 Gyr) open cluster has been comprehensively investigated by Daniel *et al.* (1994). According to Niedzielski & Muciek (1988) all selected stars are peculiar but, unfortunately, its Strömgren indices do not fall within the limits indicated for roAp stars by Martinez *et al.* (1991).

3. Observations and data reductions

The photometric observations were done with 10 or 20 sec integrations with occasional interruptions for sky background measurements. Whenever possible, we observed a comparison star in the second channel of the photometer. Even if the comparison star's data have not been directly used for the light curve of target star (due to large differences in magnitudes or of spectral types), the second channel data were adapted for estimation of the sky transparency behavior in the investigated frequency region.

Since we could obtain only short duration observations (1 to 2 hours), we developed methods for cleaning of the data obtained under not very good atmospheric conditions.

Low-frequency atmospheric and instrument trends of the data were removed by using Butterworth's filter (see Otnes & Enochson 1978). A rectangle filter was applied to comparison star data for smoothing. The programs PERIOD (Breger 1990), FOUR (Andronov 1994) and Period98 (Sperl 1998) were used for the frequency analysis. Summary of the results are given in the following section.

4. Results

From 1993 to 1997, five candidate stars of the field regions were investigated. The Fourier spectra of 4 stars namely HD 15257, HD 17317, HD 115606 and HD 217401 did not show variability in the frequency region 50–400 c/d (Dorokhova 1997). However, our persistent efforts have been rewarded with a discovery of new roAp star HD 99563 in 1997. Further details are described.

4.1 Discovery of rapid oscillations in Ap star HD 99563

HD 99563 classified as F0 in HD catalogue and F0 Sr in the Ap and Am stars catalogue by Renson (1991), its Strömgren indices ($H\beta = 2.830$, $\delta m_1 = 0.000$ and $\delta c_1 = -0.109$) fall into the limit provided by Martinez *et al.* (1991). The star belongs to a binary system ADS 8167, B component (9.^m9) is located at an angular distance of 1".7.

We observed HD 99563 during 2 hours on the night of 12th/13th May 1997 under non-photometric sky conditions with a comparison star ($\alpha(2000) = 11^h 27^m 54^s$, $\delta(2000) = -09^\circ 01'.5$) in channel 2. We presented its preliminary results at 26th Vienna Workshop on CP stars (Dorokhova & Dorokhov 1998) and Odessa "Variable stars" conference (Dorokhova 1997). In this paper, the data are reanalyzed and the dominant frequency has been refined. Fourier spectrum of HD 99563 revealed the most prominent peak (Fig. 1) at frequency $f = 130.4 \pm 0.6$ c/d (period 11 min and semi-amplitude 3.51 ± 0.3 mmag). The light curve of HD 99563 with the best-fit sinusoid of period 11 min is shown in Fig. 2.

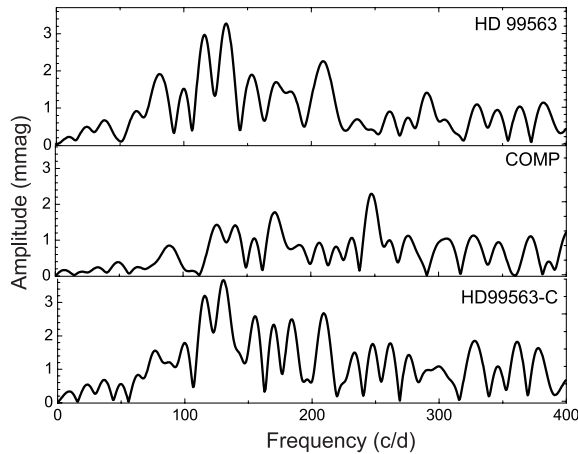


Figure 1. Fourier spectra of HD 99563 (upper), comparison star (middle) and difference (HD 99563-comp.) (bottom).

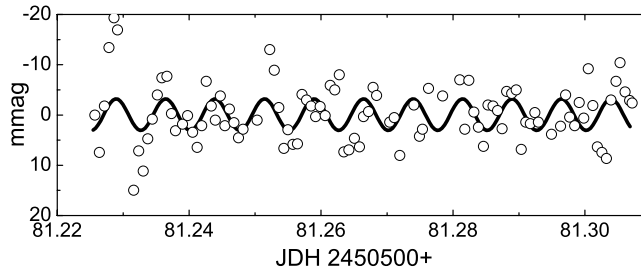


Figure 2. A light curve of HD 99563. The solid line is a least-squares fit of 11 min period.

Table 2. Journal of high-speed photometry observations during 2001–2003. All observations were obtained in *B* photometric pass band. The last two columns are duration of observations in hours and mean deviation in amplitude for the target star in the high-frequency region.

Star	Date	ΔT (hr)	σ (mmag)
HD 276625	20.11.2001	1.5	1.0
HD 281367	21.10.2001	2.2	1.6
HD 281886	24.10.2001	2.5	1.0
BD +8087	12.11.2001	2.0	3.2
	17.07.2003	5.3	1.8
	18.07.2003	5.4	1.6
	19.07.2003	2.8	3.1

A year later Handler & Paunzen (1999) had confirmed our conclusions concerning high frequency variability of HD 99563 with APT at Fairborn Observatory, Arizona. They noted that Mt. Dushak–Erekdag Observatory was the second Northern Hemisphere observatory in which roAp star had been discovered.

HD 99563 as one of the most evolved roAp stars turned out to be very interesting. In 1998, the multi-site campaign for the star revealed 8 frequencies in its light curve (Handler *et al.* 2002). In 2002, measurements of the magnetic field of HD 99563 were obtained with VLT (Hurbig *et al.* 2004). It showed unusual variations (Kurtz 2005).

4.2 Observations of 2001–2003

During 2001–2003, we observed 4 more candidates (see Table 2). All of them are classified in catalogue of Am and Ap stars (Renson 1991) as cool Ap stars with Sr, Cr or Eu. However, rapid oscillations in their data were not detected. Table 2 is a journal of observations: HD number, date, duration of observations in hours and mean deviation for high-frequency region of the target star in mmag.

We continued observing the northernmost BD + 8087 star. In a single non-ideal photometric night of 2001, target seemed to show oscillations with period of about 11 min and amplitude of 7 mmag. In the data of 2003 once more the peak at 136 c/d was revealed. However, the comparison star data also showed the peak at 142 c/d. The standard deviations of about 3 mmag also indicate the possible rapid variability although that may be due to instrumentation effects for the circumpolar stars.

HD 281886, HD 281367 and HD 276625 did not show the rapid variability during our limited observations. We plan to observe them in future also.

4.3 Negative results for stars in NGC 752 region

Four stars of NGC 752 open cluster were observed in 1993. The fourier spectra of these observations are shown in Fig. 4. R142 is a constant star within limits of 1 mmag; the frequency spectrum of R143 showed some peaks with very low amplitudes which however are not stable from night to night, and are therefore attributed to atmospheric

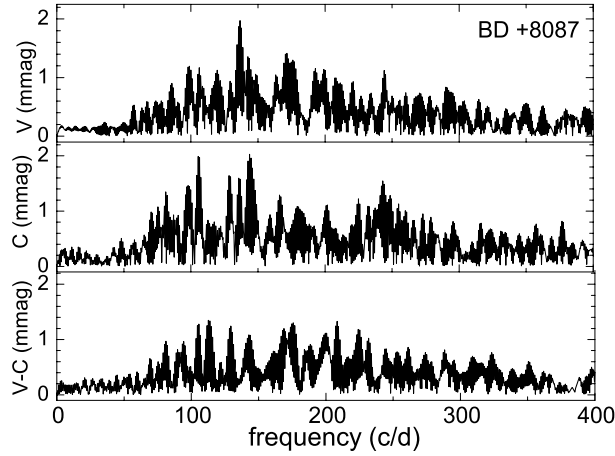


Figure 3. Fourier spectra of BD + 8087 (upper), comparison star (middle) and differences var-comp (bottom).

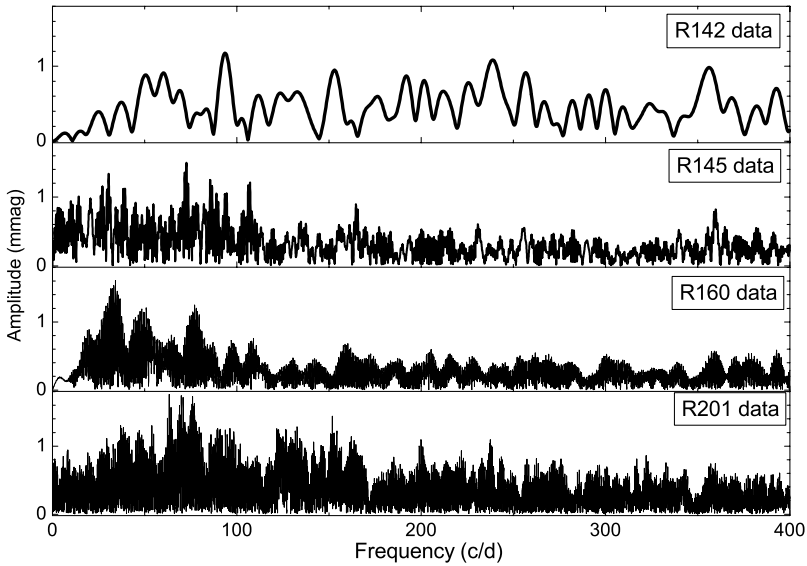


Figure 4. Fourier spectra of NGC 752 objects, the numbers in the right corners are according to Rebeiro (1970).

transparency. In the data of R160, the presence of pulsations of 34 and 77 c/d are suspected but they have very low amplitudes and are again not stable from night to night. The frequency spectrum of R201 shows some peaks in the 60–75 c/d with amplitudes of 2 mmag, which are unstable from night to night and, apparently, caused by variations of atmospheric transparency. We plan to observe them in future with CCD Camera.

5. Conclusions

Comparing with Nainital–Cape survey, which includes observations of ~ 150 stars so far (see Sagar & Mary 2005), we have observed only a little more than a dozen stars during the last 10 years. Our survey is not extensive mainly due to logistic regions. However, we have been able to discover one northern roAp star and suspect variability in a few other stars. We plan to continue our survey to search for more roAp stars.

Acknowledgements

We are very grateful to Prof. Ram Sagar, Dr. Seetha and the Local Organizing Committee for hospitality. TND expresses thanks to Prof. Don Kurtz, who was the real soul of the workshop, for his attention to our work. The data from NASA ADS, SIMBAD, CADC, and VALD databases were used.

References

- Andronov, I. L. 1994, *Odessa Astronomical Publications*, **7**, 49.
 Breger, M. 1990, *Commun. Asteroseismology*, **20**, 1.
 Daniel, S., Latham, D. W., Mathieu, R. D., Twarog, B. A. 1994, *PASP*, **106**, 281.
 Dinescu, D. I., Demarque, P., Guenther, D. B., Pinsonneault, M. H. 1995, *Astron. J.*, **109**, 2090.
 Dorokhov, N. I., Dorokhova, T. N. 1994a, *Odessa Astron. Publ.*, **7**, 167.
 Dorokhov, N. I., Dorokhova, T. N., Mkrtichian, D. E. 1994b, *Odessa Astron. Pub.*, **7**, 168.
 Dorokhova, T. N. 1997, *Odessa Astron. Publ.*, **10**, 101.
 Dorokhova, T. N., Dorokhov, N. I. 1998, In *Proc. of 26th Meeting of European Working Group on CP Stars*, p. 338.
 Girish, V., Seetha, S., Martinez, P. *et al.* 2001, *A&A*, **380**, 182.
 Handler, G., Paunzen, E. 1999, *Astron. and Astrophys. Suppl.*, **135**, 57.
 Handler, G., Weiss, W. W., Shobbrook, R. R., Garrido, R., Paunzen, E. *et al.* 2002, *ASP Conf. Proc.*, **256**, 109.
 Hauck, B., Mermilliod, M. 1990, *Astron. and Astrophys. Suppl.*, **86**, 107.
 Hauck, B., Mermilliod, M. 1998, *Astron. and Astrophys. Suppl.*, **129**, 431.
 Heller, C. H., Kramer, K. S. 1988, *PASP*, **100**, 583.
 Hubrig, S., Kurtz, D. W., Bagnulo, S. *et al.* 2004, *Astron. Astrophys.*, **415**, 661.
 Klochkova, V. G., Kopylov, I. M. 1986, *Astronomy Reports*, **63**, 240.
 Kurtz, D. W. 1990, *Ann. Rev. Astron. Astrophys.*, **28**, 607.
 Kurtz, D. W., Martinez, P. 2000, *Balt. Astr.*, **9**, 253.
 Kurtz, D. W. 2005, this volume.
 Lebedev, V. S. 1986, *AISAO*, **21**, 30.
 Milone, A. A. E., Latham, D. W. 1994, *Astr. J.*, **108**, 1828.
 Martinez, P., Kurtz, D. W., Kauffmann, G. M. 1991, *Mon. Not. R. Astr. Soc.*, **250**, 666.
 Martinez, P., Kurtz, D. W. 1994a, *Mon. Not. R. Astr. Soc.*, **271**, 118.
 Martinez, P., Kurtz, D. W. 1994b, *Mon. Not. R. Astr. Soc.*, **271**, 129.
 Matthews, J. M., Wehlau, W. H. 1985, *PASP*, **97**, 841.

- Matthews, J., Kreidl, T. J., Wehlau, W. H. 1988, *PASP*, **100**, 255.
Nelson, M. J., Kreidl, T. J. 1993, *Astron. J.*, **105**, 1903.
Niedzielski, A., Muciek, M. 1988, *Acta Astronomica*, **38**, 225.
Otnes, R. K., Enochson, L. 1978, *Applied time series analysis V.I. Basic Technologues*, New York, p. 428.
Rebeirot, E. 1970, *Astron. Astrophys.*, **4**, 404.
Renson, P. 1991, *Catalogue Genegal des etoiles Ap et Am*, p. 149.
Sagar, R., Mary 2005, this volume.
Sperl 1998, *Commun. Asteroseismology*, **111**, 1.

Using the Seismology of Non-magnetic Chemically Peculiar Stars as a Probe of Dynamical Processes in Stellar Interiors

Sylvain Turcotte

Bishop's University, Lennoxville, Canada, J1M 1Z7.

e-mail: turcotte@apollo.ubishops.ca

Abstract. Chemical composition is a good tracer of the hydrodynamical processes that occur in stars as they often lead to mixing and particle transport. By comparing abundances predicted by models and those observed in stars we can infer some constraints on those mixing processes. As pulsations in the stars are often very sensitive to chemical composition, we can use asteroseismology to probe the internal chemical composition of stars where no direct observations are possible. In this paper I focus on main sequence stars Am, λ Bootis, and HgMn stars and discuss what we can learn of mixing processes in these stars from seismology.

Key words. Variable stars—diffusion.

1. Introduction

Chemically peculiar stars (CP stars) are stars in which at least one, but typically several elements are significantly over- or under-abundant with respect to what is considered a normal composition for such stars. In population I main-sequence stars, this standard is the sun.

Chemically peculiar stars come in many forms throughout the HR diagram, and the physical mechanism(s) responsible for those anomalies vary in different regions of the HR diagram. In giant stars for example episodes of deep convection mix matter that has undergone nucleosynthesis with matter from cooler regions unaffected by nuclear processes (see Busso *et al.* 1999 for example). In cool population I and population II stars, turbulent mixing may lead to the destruction of light elements such as lithium from the surface layers of the stars (Michaud & Charbonneau 1991). The sun for example is depleted of lithium by more than two orders of magnitude. Accretion of material onto the surface of stars or mass loss by a stellar wind have also been shown as means to produce chemically anomalous stars (Proffitt & Michaud 1989; Venn & Lambert 1990). Finally, microscopic diffusion due to gravity and radiation pressure acting differently on each atomic species is another way to build abundance anomalies in some types of stars (see a recent review by Vauclair (2003) and references therein). In the following we will focus on diffusion and accretion as they are the mechanisms that are dominant in main sequence CP stars of low to intermediate mass.

The chemical anomalies observed at the surface of CP stars extend in their interior. The abundance profiles depend on the interplay of diffusion, of other processes

promoting stratification such as accretion and of mixing processes. It is these abundance profiles that we may be able to probe by studying the pulsations of those stars. If the basic physical processes leading to the formation of inhomogeneities are modeled accurately, we can then use the information on internal composition provided by pulsations to put constraints on the mixing processes that oppose the changes in composition.

In what follows we will see how chemical anomalies and pulsations are connected more specifically in main-sequence A and B-type stars where we find a large variety of both CP and variable stars. We will consider only non-magnetic stars to avoid the complications posed by the largely unknown strength and geometry of the magnetic fields in stellar interiors. Constraints on mixing processes can only be inferred when the other processes leading to the evolution of the composition can be modeled with no free parameters. We will concentrate on Am and HgMn stars, main sequence A and B type stars respectively, in which diffusion is the governing process and on λ Bootis stars that are thought to be formed through accretion of metal depleted circumstellar matter.

2. Abundance stratification in main-sequence stars

The governing equation for the evolution of abundances in a star is

$$\frac{\partial X}{\partial t} = \left[(D_i + D_T) \frac{\partial X}{\partial M} + (v_i + v_w + v_{\text{hydro}}) X \right] - \lambda_{\text{sink}} X + \lambda_{\text{source}}, \quad (1)$$

where X is the mass fraction of a species. D_i and v_i are the diffusion coefficient and velocity for that species, D_T is the turbulent diffusion coefficient, v_w is a velocity due to mass loss (if positive) or accretion (if negative), v_{hydro} is a velocity due to large scale motions of matter (meridional circulation f. ex.), λ_{sink} is a sink term (nonzero at the surface in the case of mass loss), and λ_{source} is a source term (nonzero at the surface in the case of accretion).

The important processes that lead to abundance stratification (and surface abundance anomalies) and the competing mixing mechanisms contribute differently in each star depending on its temperature, age, mass, and rate of rotation. Accretion on the other hand, when present, is an external process practically independent of the properties of the star.

2.1 Mixing

Several processes lead to mixing in stars, in some cases through large scale motion that transport matter throughout the interior, in other cases through local motions such as turbulence. The most common and most important of the mixing mechanisms is convection. In so-called *standard* stellar models, convection is typically the only mixing mechanism considered. Amongst the other mixing processes the most important are convection overshoot and rotational induced mixing. Rotation leads to not only large scale motions, meridional circulation but also to turbulent motions induced by rotational shear or as the result of the transfer of angular momentum from the core toward the surface. In stellar models turbulent mixing due to convection or other forms of turbulent motions is accounted for through the turbulent diffusion coefficient (D_T in equation 1). In the models discussed below, D_T for convection is arbitrary but must

be very large, while the functional form for the combined effect of all other mixing mechanisms is an *ad hoc* power law (Richer *et al.* 2000).

Convection typically occurs in regions of partial ionization of HI and He I at around 10 000 K and of He II at around 40 000 K. In non-magnetic stars with surface temperature of less than 10 000 K, the outer regions of the star, including the photosphere, are convective. The HI + He I and He II convection zones are expected to be linked by convection overshoot. As a result, the whole outer envelope of these stars are completely mixed. Mixing is also expected to extend below the He II convection zone because of overshoot or turbulence. The region in which the star is fully mixed extends to a certain depth. This mixed region is named the Superficial Mixed Zone (SMZ) in the remainder of this paper. The extent of the SMZ may be determined by seismology.

In cool stars, the photospheric abundance measured spectroscopically are identical to the abundances at the deepest point of the SMZ. Surface abundances are determined by particle processes that occur relatively deep in the interior. In stars hotter than 10 000 K, the outermost regions are not convective and therefore may not be mixed. Stratification can therefore occur in the photosphere. As a result, the photospheric abundances can be completely different from the abundances in the star's interior.

2.2 Diffusion

Chemical species drift relative to each other as the net forces acting on each species vary depending on their atomic properties. Essentially, it is the competition between gravity and radiation pressure that determines if an element will levitate toward the surface or sink toward the center of the star. While gravity is essentially constant in the outer parts of stars, radiation pressure goes through peaks and troughs depending on the local temperature and density. Where radiation pressure is larger than gravity, an element will drift outward, but will sink if gravity is larger. Figure 1 shows a cartoon that illustrates the effect of diffusion on the chemical profiles and the effect of the depth of the SMZ on surface abundances. In the middle panel, the base of the SMZ occurs where gravity is larger than radiation pressure so the abundances in the SMZ are lower than initially. In the bottom panel, the deeper SMZ yields an increase in the surface composition but a smaller anomaly due to the larger mass of the SMZ.

The effect of diffusion on abundances can be calculated rather accurately from first principles. As a result, stars in which diffusion is the dominant process in creating abundance gradients are very useful tools to study mixing processes.

2.3 Accretion

Accretion does not depend directly on intrinsic properties of the accreting star but that the material falling on the surface of the star must be mixed (diluted) in the entire SMZ. The composition in the mixed zone evolves on a time scale defined by the ratio of the mass of the mixed zone to the accretion rate ($\tau = M_{\text{SMZ}}/\dot{M}$; neglecting here the effect of other processes at the base of the mixed zone). The depth in the star that the accreted material will reach depends on the depth of the SMZ, but also on the advection of the accreted matter in the star at a velocity of $v = \dot{M}/4\pi\rho r^2$ (Charbonneau 1991).

Given a large enough rate of accretion (*i.e.*, a short τ compared to the evolution timescale), the abundances in the SMZ will be identical to those of the accreted matter. Once the accretion episode ends, the signature of accretion rapidly disappears (see

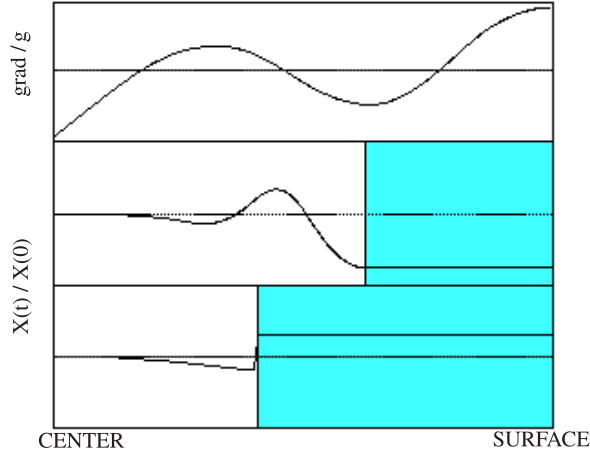


Figure 1. The top panel is a cartoon representation of the ratio between the acceleration due to radiative pressure (g_{rad}) and the gravitational acceleration (g), where the horizontal dotted line shows a ratio of one. The middle panel shows the evolution of the abundance profile as the result of the g_{rad}/g profile shown above. The shaded region on the right is the Superficial Mixed Zone (SMZ) in which the abundance is homogenized by mixing. The bottom panel shows the evolution of the same element but for a deeper SMZ.

detailed models in Turcotte & Charbonneau 1993). The abundance profiles will depend on the SMZ, on $\dot{M}$, and on the cumulative mass of the accreted material. None of which can be measured directly. As a result it may be difficult to find the depth of the SMZ uniquely for a given star.

3. Composition anomalies and their effect on pulsations

Different elements dominate the opacity at different depths in a star. Hydrogen and helium contribute most of the opacity at low temperature, iron dominates in hotter regions. In CP stars, the opacity profile will be different compared to chemically homogeneous stars. The opacity might be smaller if the elements that dominate at a given temperature become less abundant, and inversely, might increase elsewhere as some elements accumulate where they play an important role.

In Fig. 2, the contribution of several elements to the opacity is shown in the case of a chemically homogeneous star of solar composition and in a star in which the abundances have been changed by diffusion. One can notice that the contributions of helium and of CNO have been reduced in the outer regions of the star while the contributions of iron-peak elements have increased. As a result, the driving due to helium is smaller in stars with diffusion than in normal stars, while the driving due to heavy elements is enhanced.

In stars in which accretion is ongoing, the effect on driving will be determined by the content of the accreted material and the depth in the star reached by the accreted material. In λ Bootis stars, the driving due to helium will be unchanged as helium is expected to be present in the accreted material and the helium driving region is inside the surface convection zone. The driving due to heavy elements will be changed only in certain circumstances.

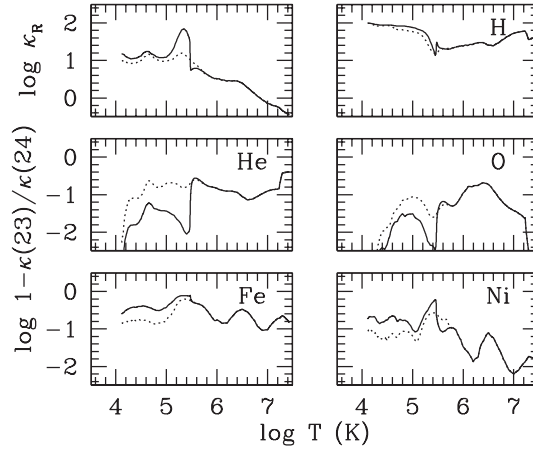


Figure 2. The relative contribution of individual elements to the Rosseland mean opacity is shown for 5 elements when the chemical composition is uniformly solar (dotted line) and when diffusion has occurred (solid line). The top left panel shows the logarithm of the Rosseland mean opacity in both cases. In the five other panels the logarithm of the ratio of the opacity including all elements in the star ($\kappa(24)$) and the opacity calculated without the contribution of the element labeled ($\kappa(23)$) is plotted. If an element provides all the opacity, then the curve for that element would reach one. The model shown is a $4 M_{\odot}$ star model of 100 Myr (see Turcotte & Richard 2005, submitted).

4. Inference of mixing processes in CP stars

The ideas above have been applied to three types of chemically peculiar stars of the main sequence, the A-type Am and λ Bootis stars, and the B-type HgMn stars. Although similar in many ways these three types of stars are nonetheless different in important ways: when variable, Am and λ Bootis stars are δ Scuti type pulsators (p -modes driven by helium) while pulsating late-type B stars are g -modes excited by the opacity of heavy elements; Am and HgMn stars are slowly rotating while λ Bootis stars are moderately rapid rotators; and, finally, because of differences in the position of convection zones, the photospheric composition of Am and λ Bootis stars is the same as that of the driving region while it is not the case for HgMn stars because of diffusion in their atmosphere. The expected differences among these three types of stars are illustrated in Fig. 3.

4.1 Application to Am stars

Models of non-rotating A stars were computed to test whether variable Am stars can be driven by the opacity of helium as in δ Scuti stars (Turcotte *et al.* 2000). As suggested in Fig. 2, when the helium settles out of the driving region as a result of diffusion the driving in Am stars is much lower than in chemically normal stars. As the helium driving region is inside the superficial convection, and thus inside the SMZ, the level of depletion of helium in the driving region is a direct function of the depth at which the SMZ extends; the deeper the SMZ, the higher the driving will be.

The blue (hot) edge of the instability zone for Am stars is in direct relation to the abundance of helium in the driving region as is shown in Fig. 4. Therefore, we can infer the depth of mixing from the instability region for Am stars. The boundary of the zone

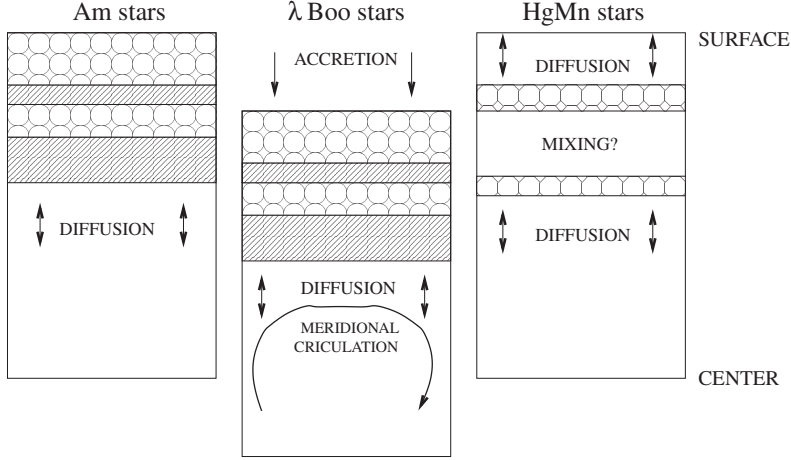


Figure 3. The cartoons (not to scale) illustrate the expected processes important in the evolution of the abundances of the three types of stars discussed in the paper. The regions filled with circles represent convection zones, the regions filled by oblique lines are non-convective regions but otherwise mixed by other turbulent processes. The presence or absence of non-convective mixed regions in HgMn stars is not known. The deeper convection zone in HgMn stars appears only when iron-peak elements have accumulated in that region as the result of diffusion.

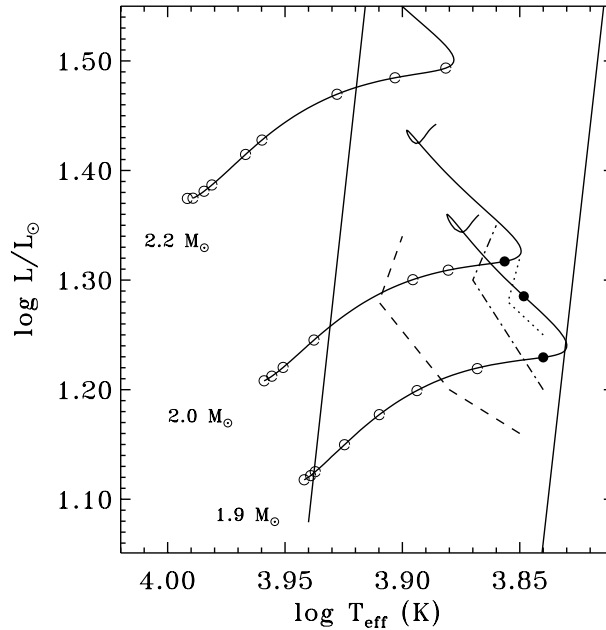


Figure 4. The HR diagram shows the approximate boundaries of the classical instability strip (oblique lines), the evolutionary paths of 1.9, 2, and 2.2 $M_{\odot}$ stars undergoing diffusion. The filled circles represent models where at least one unstable pulsation mode was found, the open circles show models that are stable. The three other lines show the approximate boundaries of the instability region for variable Am stars when the SMZ is shallow (dotted line), best fit (dash-dotted), or deep (dashed).

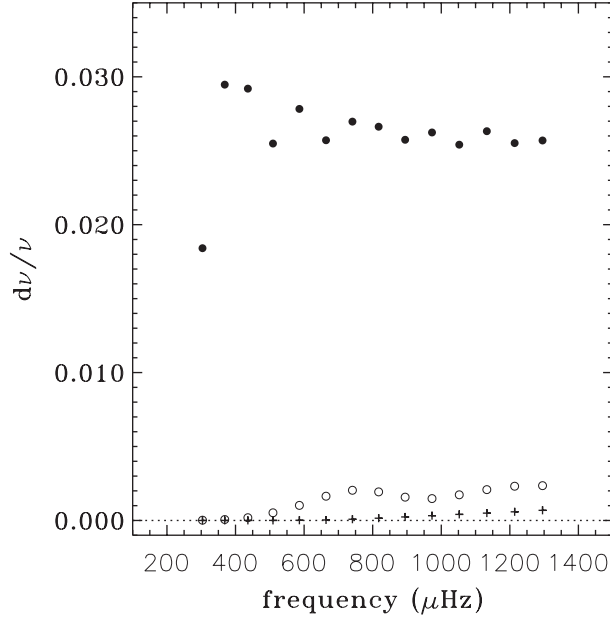


Figure 5. The relative difference in frequencies of radial modes of order 1 to 14 for models that are metal poor (except for CNO and S) to a depth where the temperature is $\log T = 4.4$ (crosses), $\log T = 5.0$ (open circles), and $\log T = 5.8$ (filled circles) with respect of a chemically homogeneous model are shown at the frequency of the reference model. The SMZ in the last model is 10^6 more massive than in the first.

has not been explored in detail yet, partly because of the computational cost to do so. An additional difficulty is that the initial helium content of individual Am stars need not be identical and so the depth of mixing may not be a unique solution for all variable Am stars. Nevertheless, the depth of mixing found through seismology is consistent with what was found from modeling the surface composition (Richer *et al.* 2000).

4.2 Mixing and accretion in λ Bootis stars

In the accretion model the light elements (H, He, CNO) in the SMZ are replenished by the infalling gas while most other elements are depleted by a factor of 10 in the accreting matter (Turcotte & Charbonneau 1993). As a result, one would not expect to see any measurable effect on the pulsations of λ Bootis stars in comparison to δ Scuti stars if the SMZ is not significantly deeper than the surface convection zone of normal A stars. If, however, the SMZ extends significantly deeper, to the depth at which the opacity of heavy elements drive pulsations, an observable signature might be generated. Figure 5 shows the relative change in frequency for radial modes ($l = 0$) of radial order (n) of 1 to 14 with respect to a chemically homogeneous model. We see that the differences are significant only if the chemical anomalies extend to a depth where the low metallicity of the accreted matter can make a large difference in the opacity ($\log T > 5.2$). Models of Am stars suggest this should be the case, and as λ Bootis stars are faster rotators one may expect more mixing in λ Bootis stars than in Am stars.

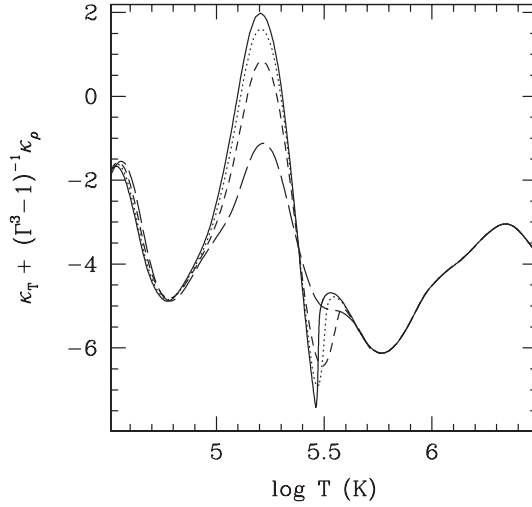


Figure 6. The sum of the logarithmic derivatives of opacity ($\kappa_T = d \log \kappa / d \log T$, $\kappa_\rho = d \log \kappa / d \log \rho$) is shown for four $4 M_\odot$ models with more or less mixing versus the temperature in the star. The peak at around $\log T = 5.2$ is due to the opacity of iron-peak elements. It is larger in models where mixing is lower and abundance anomalies are allowed to be larger. The pulsations are driven in regions where $d/dr[\kappa_T + \kappa_\rho/(\Gamma_3 - 1)] > 0$.

Unfortunately, using frequencies to learn about the structure of δ Scuti type stars are yet to be fruitful due to problems in mode identification and in dealing with rotation theoretically. Consequently, relying on the small frequency differences between the different models to infer the depth of the SMZ promises to be challenging.

4.3 What is the matter with HgMn stars?

In pulsating B stars the driving is due to iron-peak elements (Pamyatnykh 1999). As HgMn stars are slowly rotating, the current theoretical expectation is that diffusion will occur in the interior as it does in the atmosphere (as the large abundance anomalies measured there attest). Models show that diffusion leads to an accumulation of iron-peak elements in the driving region. The increase in abundance of those elements naturally leads to an increase in opacity, which in turn leads to an enhanced driving. Figure 6 shows the logarithmic opacity gradients for models of a $4 M_\odot$ where the depth of the SMZ is varied. The increase of the peak at $\log T \approx 5.2$ leads to an increase in the derivative of these opacity gradients which means that the driving does increase (see Shibahashi in these proceedings).

However, HgMn stars are not known to be variable whereas chemically normal stars of the same mass and evolutionary state are variable (those are the Slowly Pulsating B stars). The disagreement between the models and observations suggests that the models are deficient in some ways. A solution may come from introducing new physics into the models such as selective mass loss, but another possibility may simply be that the models are not accurate enough in the superficial regions. Indeed, the models discussed here assume for reasons of numerical stability that the stars are fully mixed from the photosphere to a depth past the mode driving region which is not realistic (Turcotte & Richard 2005, submitted). Only through further modeling combined with

seismic data will we establish whether HgMn stars are really stable or variable at very low amplitude and what implication, if any, this will have on dynamical processes occurring in these stars.

5. Conclusions

Using variability to infer mixing is especially useful in stars for which surface abundance is not a good indicator of internal composition, such as HgMn stars, and, to a certain degree, λ Bootis stars. In Am stars on the other hand, variability is an indicator of the abundance of helium, which is useful as it cannot be seen in the spectra of these stars.

Much more work is required to refine this tool. At this time only the depth of the superficial region where abundances are homogeneous (the SMZ) can be constrained in Am stars. The exact processes that lead to that mixing cannot be isolated nor can the exact profile of the turbulent diffusion coefficient be determined. Nevertheless, this tool remains our best hope to study the internal chemical composition of stars of the upper main-sequence. As models improve the data from MOST, CoRot and other missions will hopefully provide us with low amplitude modes in Am, λ Bootis, and HgMn stars.

References

- Busso, M., Gallino, R., Wasserburg, G. J. 1991, *Ann. Rev. Astron. Astroph.*, **37**, 239.
Charbonneau, P. 1991, *Astrophys. J. Letters*, **372**, 33.
Michaud, G., Charbonneau, P. 1991, *Space Sci. Rev.*, **57**, 1.
Pamyatnykh, A. A. 1999, *Acta Astronomica*, **49**, 119.
Proffitt, C. R., Michaud, G. 1989, *Astrophys. J.*, **345**, 998.
Richer, J., Michaud, G., Turcotte, S. 2000, *Astrophys. J.*, **529**, 338.
Turcotte, S., Charbonneau, P. 1993, *Astrophys. J.*, **413**, 376.
Turcotte, S., Richer, J., Michaud, G., Christensen-Dalsgaard, J. 2000, *Astron. Astrophys.*, **360**, 603.
Vauclair, S. 2003, *Astrophysics and Spa. Sci.*, **284**, 205.
Venn, K. A., Lambert, D. L. 1990, *Astrophys. J.*, **363**, 234.

Asteroseismology of δ Scuti and γ Doradus Stars

Gerald Handler

*Institut für Astronomie, Universität Wien, Türkenschanzstrasse 17, A 1180 Wien, Austria.
e-mail: handler@astro.univie.ac.at*

Abstract. We give an overview of past and present efforts to make seismology of δ Scuti and γ Doradus stars possible. Previous work has not led to the observational detection and identification of a sufficient number of pulsation modes for these pulsators for the construction of unique seismic models. However, recent efforts including large ground-based observational campaigns, work on pre-main sequence pulsators, asteroseismic satellite missions, theoretical advances on mode identification methods, and the discovery of a star showing simultaneous self-excited δ Scuti and γ Doradus oscillations suggest that we may be able to explore the interiors of these pulsators in the very near future.

Key words. Stars: variables: δ Scuti—stars: oscillations, variables: other.

1. Introduction

The δ Scuti and γ Doradus stars are pulsating variables of spectral types A or F that are located on or near the main sequence. Although they occupy similar regions in the HR diagram (see Fig. 1), there are a number of important differences between the pulsational behaviours of these two groups of variable stars.

For instance, the δ Scuti stars pulsate with periods between about 20 minutes and 8 hours. Therefore they oscillate in pressure (p) and gravity (g) modes of low radial order. On the other hand, the pulsation periods of the γ Doradus stars range from 8 hours to 3 days, indicating that g modes of high radial order are excited. Although there is some overlap in the pulsation periods of the two classes of pulsators, they are clearly separated if the pulsation “constant” Q is considered (Handler & Shobbrook 2002).

Both the δ Scuti and γ Doradus stars contain variables that oscillate in multiple modes. This makes them interesting for asteroseismic studies and several attempts have been made to explore their interior structures by modelling their pulsation spectra. However, these attempts have not been particularly successful to date.

Consequently, the remainder of this paper is intended to give an overview of the attempts to apply asteroseismology to δ Scuti and γ Doradus stars. We first identify the areas in which problems for asteroseismic studies have been met, and then try to show how they can be overcome.

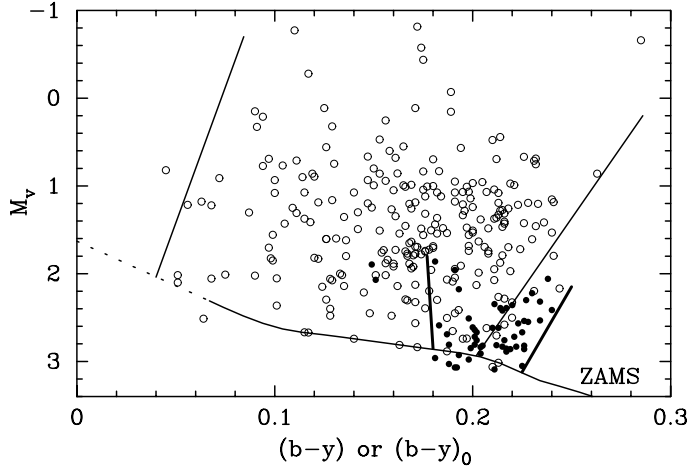


Figure 1. The location of δ Scuti and γ Doradus stars in a colour-magnitude diagram. The open circles represent all δ Scuti stars for which Strömgren colours and HIPPARCOS parallaxes are available, whereas the filled circles show all γ Doradus stars with the same information. The Zero-Age Main Sequence (ZAMS) is indicated, taken from the standard relations by Crawford (1975, 1979), and is extrapolated towards low $(b-y)$ (dotted part of the ZAMS). The observational borders of the two pulsational instability strips (Rodríguez & Breger 2001; Handler & Shobbrook 2002), with the thicker lines corresponding to the γ Doradus strip, are also shown.

2. Challenges for asteroseismology

The two main interior regions of a δ Scuti star comprise a convective core and a (mostly) radiative envelope. Both of these regions contain a specific family of pulsation modes: gravity modes are in principle located in the core, whereas pressure modes mainly propagate in the envelope. At the ZAMS, the periods of the g modes are considerably longer than those of the p modes and the two sets of mode spectra are well separated in period.

However, as a δ Scuti star evolves along the main sequence, its interior structure changes. The convective core shrinks and becomes more compact, whereas the envelope expands and its mean density decreases. This means that the g -mode periods decrease, and at the same time the p -mode periods increase. At some point, the period ranges of these two mode families begin to overlap, modes start to exchange properties (resulting in so-called mixed modes), and the avoided crossing phenomenon (e.g., see Aizenman *et al.* 1977) takes place.

This has a dramatic effect on the theoretically predicted mode spectra. Whereas a pure p -mode spectrum should be observable at the ZAMS, an ever-increasing number of mixed modes is predicted to be excited during the late- and post main sequence stages of stellar evolution (e.g., see Dziembowski & Krolikowska 1990). We illustrate this problem in Fig. 2.

The number of theoretically predicted pulsation modes increases considerably as a model reaches the end of its main sequence life, and the theoretical pulsation spectrum of post main sequence models is extremely dense. For comparison, model A has a total of 24 unstable pulsation modes (8 of them being $m = 0$), model B has 80 unstable modes (24 of $m = 0$), and model C has 345 unstable modes (89 of which are $m = 0$).

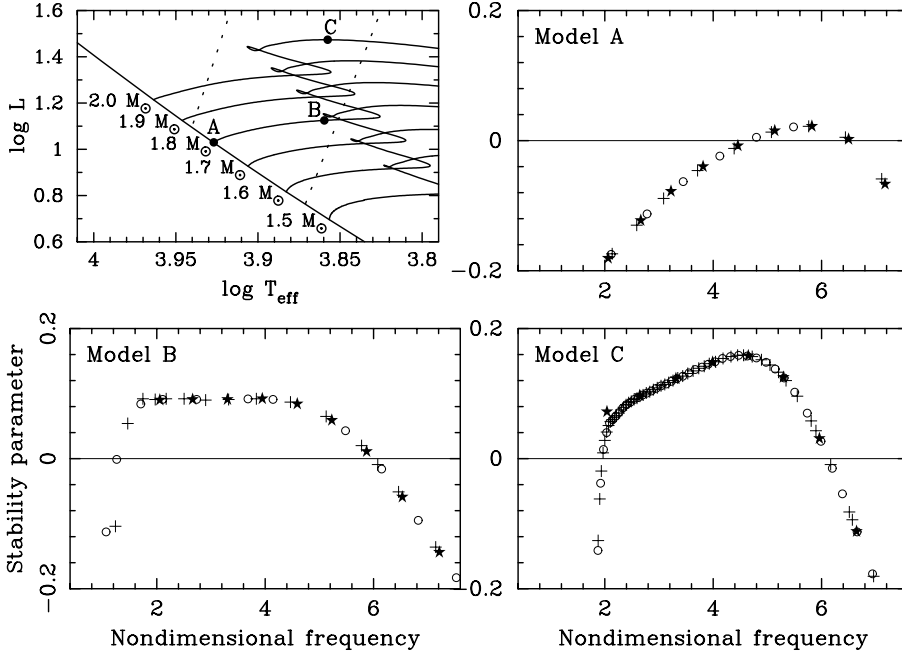


Figure 2. Upper left: a theoretical HR diagram, with some evolutionary tracks (labelled with their masses) indicated. The dotted lines are the borders of the δ Scuti instability strip. Three selected models in different evolutionary stages (ZAMS, model A; end of main sequence, model B, post main sequence, model C) are indicated. Upper right and lower two panels: pulsationally stable and unstable modes (only the $m = 0$ components are plotted) in models A, B and C. All modes with a positive stability parameter are excited in the models. Star signs indicate radial modes, open circles are $l = 1$ modes and plus signs correspond to $l = 2$ modes.

Many δ Scuti stars are also fast rotators. This means that the frequency splittings due to rotation may become large enough that their size becomes comparable to the frequency difference of consecutive radial overtones and mode coupling, leading to additional modifications of the pulsation frequencies, may also occur (see Pamyatnykh 2003).

Since ground-based observational efforts to determine the multiple pulsation frequencies of δ Scuti stars have so far revealed only a few tens of independent pulsation modes, it thus seems logical to observe unevolved pulsators that do not rotate rapidly to maximise the ratio between the number of observed and predicted pulsation modes. Such a strategy may lead to unambiguous mode identifications and hence enable seismic modelling.

Handler *et al.* (2000) followed this approach and determined a total of 22 independent pulsational mode frequencies for the δ Scuti star XX Pyxidis. This object is unevolved and about two thirds of the theoretically predicted $l \leq 2$ modes should have been detected. Still, all attempts to derive a mode identification were in vain, as no clear patterns within the mode frequencies could be detected. This example makes it clear that one needs mode identification methods to understand the pulsation spectra of δ Scuti stars.

Several of these methods exist, including photometric and spectroscopic techniques, and combinations of these (e.g., see Reed 2003 and Telting 2003 for reviews).

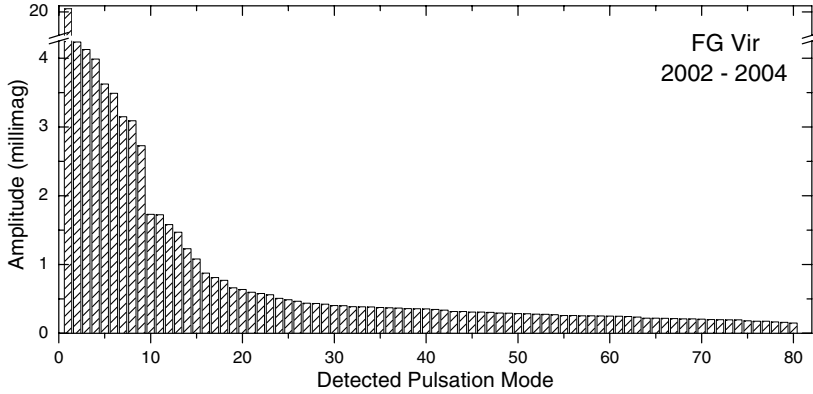


Figure 3. The number of frequencies found in the light curves of the δ Scuti star FG Vir depending on the detection threshold. Used with permission from Breger *et al.* (2005).

Unfortunately, their application also seems to be problematic for δ Scuti stars (e.g., Balona 2000), because some apparently independent methods are redundant and because the shallow surface convection zones of these stars seem to make mode discrimination ambiguous.

Summarising, we can state that the two major problems for seismology of δ Scuti stars are the detection of a sufficient number of pulsation modes and their unambiguous identification with the pulsational quantum numbers k , the radial order, l , the spherical degree and m , the azimuthal order. These difficulties must be overcome before detailed theoretical seismic studies can commence.

Turning now briefly to the γ Doradus stars, we face a similar situation. In fact, it is even more difficult to detect and identify the individual pulsation modes of these stars, since their pulsation periods are of the same order as their rotation periods, leading to considerable overlap of neighbouring mode multiplets, and complicating the theoretical treatment of rotation. In addition, the longer pulsation periods themselves require even larger quantities of data to be obtained with the most stable equipment.

Taking the discussion above at face value, it seems that asteroseismology of δ Scuti and γ Doradus stars is extremely difficult from the observational point of view. In fact, several researchers (including the present author) have stopped their work on these problems. However, such an attitude may be too pessimistic.

3. Light at the end of the tunnel?

3.1 δ Scuti stars

Fortunately, the continued efforts of several groups of researchers, including the advent of asteroseismic satellite missions have brought some important advances in both observation and theory. For instance, Breger *et al.* (2005) carried out extensive multisite ground-based photometric observations of the star FG Vir over several consecutive seasons, leading to the detection of a total of 67 independent pulsation modes (see Fig. 3).

The position of FG Vir in the HR diagram is close to that of our model B from the previous section that had a total of 80 unstable pulsation modes. Consequently, it is possible that a sufficient number of modes has been detected for seismic modelling of this star, unless a significant fraction of the observed modes is $l > 2$.

The study of pre-main sequence (PMS) δ Scuti stars also bears some seismic promise. Since the interior structure of these objects (e.g., Suran *et al.* 2001) is presumed to be relatively simple, it may also be possible to decipher their pulsational mode spectra more easily. Recent searches for PMS pulsators have brought the number of known objects up to two dozen (K. Zwintz, private communication), increasing the chances to find promising targets considerably. In fact, multisite observations of PMS δ Scuti stars have already revealed as many as 19 pulsation frequencies (T. Kallinger, private communication) for an individual star.

The biggest advances in observational asteroseismology of δ Scuti and γ Doradus stars can however be expected from present and future satellite missions. The high measurement accuracy from space will allow the detection of an unprecedented number of pulsation modes, as can also be expected from Fig. 3 where the number of detected oscillations increases sharply with decreasing detection limit.

The MOST (Microvariability and Oscillations of STars, Walker *et al.* 2003) satellite is already delivering data, with excellent performance (Matthews, these proceedings). The COROT (Convection, Rotation, and planetary Transits, e.g., Baglin 2003) mission is due to be launched in mid 2006. It contains a larger telescope than MOST (27 vs. 15 cm aperture), will monitor its targets over a longer time baseline (~ 150 d vs. ~ 40 d), and has several δ Scuti stars as primary targets. Consequently, it can be expected that COROT will deliver even better data after a successful launch.

However, it should be clearly pointed out that the advent of space asteroseismology will not make ground-based observations unnecessary. First, the satellites are strongly limited in terms of target brightness and observable region in the sky. Second, the length of the data sets obtained from space is usually not sufficient to resolve the frequencies of all the possibly excited pulsation modes in, e.g., δ Scuti stars (see Handler 2004 for a discussion). Third, the use of mode identification methods is not implemented in the present asteroseismology satellite setups. The larger flexibility of ground-based measurements will thus still make them an asset for future asteroseismic observations.

Having mentioned mode identification methods, it is noted that considerable progress has been made on the theoretical side concerning δ Scuti stars. Daszyńska-Daszkiewicz *et al.* (2003) showed that if time-resolved multicolour photometry of such a pulsator in at least three filters is available, it can be used both to infer the type of the pulsation mode as well as to obtain constraints on surface convection.

3.2 γ Doradus stars

As opposed to the δ Scuti stars, the γ Doradus stars are known to be pulsating variables only for a few years (Kaye *et al.* 1999), and hence they are not that well investigated. On the other hand, their long pulsation periods make them well suitable for measurements with automated photometric telescopes. Such searches (most recently Henry *et al.* 2005) have increased their number to a total of 54, four times as many as were known only five years ago!

Figure 4 shows a colour-magnitude diagram of the γ Doradus stars, where it becomes clear that about half of them are located within the δ Scuti instability strip as well. There

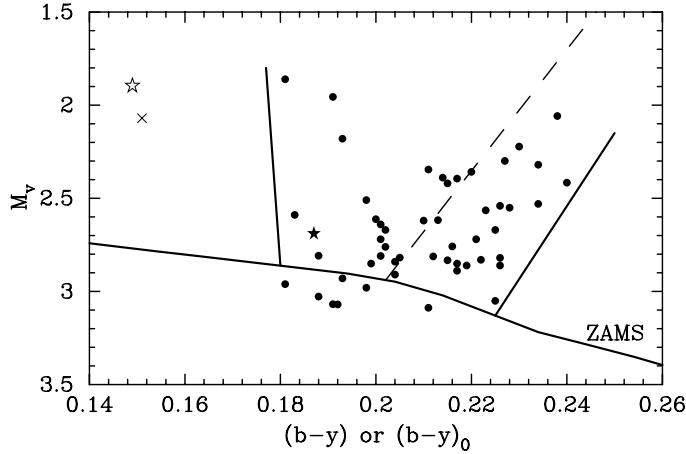


Figure 4. A colour-magnitude diagram of the bona fide γ Doradus stars (filled circles), similar to Fig. 1. The dashed line is the cool border of the δ Scuti instability strip. The open star symbol corresponds to HD 209295, whose γ Doradus pulsations are likely to be tidally excited, the cross is for HD 221866, a binary with a γ Doradus component, and the filled star symbol is HD 8801, both a γ Doradus and a δ Scuti star.

are also two prominent outliers to the blue edge of the γ Doradus region: HD 209295 is a 3-day binary in an eccentric orbit with an unknown companion that exhibits both δ Scuti and γ Doradus pulsations, but the latter are probably excited by tidal interaction (Handler *et al.* 2002). HD 221866 is also a binary, consisting of an Am star and an F2 dwarf (Kaye *et al.* 2004), and it is therefore suspected that the pulsations originate from the F2 star whereas the system’s colours are dominated by the earlier-type companion.

In any case, the stars that are located in both instability domains are particularly interesting for asteroseismology as they may show self-excited pulsations of both types. Unfortunately, searches for such stars (Handler & Shobbrook 2002) were unsuccessful—until recently. Henry & Fekel (2005) reported the discovery of simultaneous δ Scuti and γ Doradus pulsations in the apparently single Am star HD 8801. Although only a few pulsation modes were detected for this object, it is important to know that both types of pulsation can exist in the same star. It is hoped that more such objects will be discovered in the near future, since they offer the possibility to study their innermost regions, and to sound their envelopes at the same time.

In this context it is interesting to note that the γ Doradus stars (as well as the cooler δ Scuti stars) occupy a region in the HR diagram where solar-like oscillations are also predicted (Zerbi *et al.* 2000) and high-precision radial velocity observations as well as space photometry could reveal both types of variability. This would be particularly exciting as one of the unsolved questions of helioseismology is just the reverse: the presence of g modes in the sun.

4. Summary and conclusions

To date, it has not been possible to perform detailed asteroseismic studies of δ Scuti and γ Doradus pulsators. From the observational point of view, the main problems to be overcome are the detection of a sufficient number of pulsation modes, and their unique

match with the pulsational quantum numbers k , l and m . The γ Doradus stars pose additional complications due to their longer pulsational time scales that make observations and mode identifications even more difficult and that also make the theoretical treatment of rotation more elaborate.

However, some recent studies of δ Scuti stars turned out to be quite encouraging: large numbers of pulsation modes are detectable from the ground if the stars are observed persistently, pre-main sequence objects may have pulsation spectra simple enough for mode identification via pattern recognition, and some mode identification methods have matured to a point where they may be applied reliably and could even yield additional constraints on surface convection for stellar modelling. High-precision space observations will soon be available for a number of objects and it is reasonable to expect that a significantly larger number of pulsation modes will be detected as compared to ground-based measurements (which can however not be replaced by the satellite data).

There is also considerable progress in the study of the γ Doradus pulsators: we know now more than 50 such objects, and the first star that shows both self-excited γ Doradus and δ Scuti oscillations has been discovered. In addition, it also seems possible that solar-like oscillations can be present in some γ Doradus and some cooler δ Scuti stars.

Acknowledgements

I would like to thank the organisers of this meeting for their invitation and a travel grant, and in particular for their warm hospitality that made this conference an unforgettable occasion.

References

- Aizenman, M., Smeyers, P., Weigert, A. 1977, *A&A*, **58**, 41.
 Baglin, A. 2003, *Adv. Sp. Res.*, **31**, 345.
 Balona, L. A. 2000, *MNRAS*, **319**, 606.
 Breger, M. *et al.* 2005, *A&A*, in press.
 Crawford, D. L. 1975, *AJ*, **80**, 955.
 Crawford, D. L. 1979, *AJ*, **84**, 1858.
 Daszyńska-Daszkiewicz, J., Dziembowski, W. A., Pamyatnykh, A. A. 2003, *A&A*, **407**, 999.
 Dziembowski, W. A., Krolukowska, M. 1990, *Acta Astr.*, **40**, 19.
 Handler, G. *et al.* 2000, *MNRAS*, **318**, 511.
 Handler, G., Shobbrook, R. R. 2002, *MNRAS*, **333**, 251.
 Handler, G. *et al.* 2002, *MNRAS*, **333**, 262.
 Handler, G. 2004, In *Second Eddington Workshop: Stellar structure and habitable planet finding*, ESA SP-538, p. 127.
 Henry, G. W., Fekel, F. C. 2005, *AJ*, **129**, 2026.
 Henry, G. W., Fekel, F. C., Henry, S. M. 2005, *AJ*, **129**, 2815.
 Kaye, A. B., Handler, G., Krisciunas, K., Poretti, E., Zerbi, F. M. 1999, *PASP*, **111**, 840.
 Kaye, A. B., Gray, R. O., Griffin, R. F. 2004, *PASP*, **116**, 558.
 Pamyatnykh, A. A. 2003, *Ap&SS*, **284**, 97.
 Reed, M. D. 2003, *Ap&SS*, **284**, 73.
 Rodriguez, E., Breger, M. 2001, *A&A*, **366**, 178.
 Suran, M., Goupil, M.-J., Baglin, A., Lebreton, Y., Catala, C. 2001, *A&A*, **372**, 233.
 Telting, J. H. 2003, *Ap&SS*, **284**, 85.
 Walker, G. *et al.* 2003, *PASP*, **115**, 1023.
 Zerbi, F. M., Aerts, C., Handler, G., Kaye, A. B., Poretti, E. 2000, In *The Third MONS Workshop: Science Preparation and Target Selection*, Aarhus University, p. 91.

Asteroseismology of δ Scuti Stars: Problems and Prospects

M.-J. Goupil¹, M. A. Dupret¹, R. Samadi¹, T. Boehm², E. Alecian¹,
J. C. Suarez³, Y. Lebreton⁴ & C. Catala¹

¹ LESIA, UMR 8109, Observatoire de Paris, 5 place J. Janssen, 92195 Meudon, France.

² Observatoire de Midi-Pyrénées, Toulouse, France.

³ Instituto de Astrofísica de Andalucía, Granada, Spain.

⁴ GEPI, UMR 8111, Observatoire de Paris, 5 place J. Janssen, 92195 Meudon, France.

Abstract. We briefly outline the state-of-the-art seismology of δ Scuti stars from a theoretical point of view: why is it so difficult a task? The recent theoretical advances in the field that these difficulties have influenced are also discussed.

Key words. Oscillations— δ Scuti stars—stars: main and pre-main sequence stars.

1. Introduction

This review is intended for students and readers who are unfamiliar with δ Scuti stars. For more detailed and specific reviews on this topic, we refer to Dziembowski (1990); Goupil & Talon (2002); Pamyatnykh (2003); Kurtz (2003); Poretti *et al.* (2003); Breger *et al.* (2004).

Section 2 defines what is meant here by a δ Scuti star. Section 3 explains the main problems which are responsible for the fact that the seismology of δ Scuti stars has not yet held its promises. Section 4 presents some theoretical advances which can help to overcome some of the above difficulties. Section 5 introduces a field which is rapidly developing: seismology of oscillating pre-main sequence (PMS) stars which are located roughly at the same location in the HR diagram but are moving on their way towards the main sequence instead of evolving on or off the main sequence (reviews Catala 2003; Marconi & Palla 2004).

2. What is a δ Scuti star?

δ Scuti stars are intrinsically pulsating stars of spectral type A–F, with masses ranging roughly between 1.5 and 2.5 $M_{\odot}$. They evolve on or near the main sequence in the HR diagram and cross the so-called classical (or Cepheid) instability strip.

2.1 Structure and evolution

The classical δ Scuti stars lie on the main sequence where they burn their central hydrogen. These stars have a convective core which recedes with time leaving behind

a chemical gradient. The temperature gradient is so large that the nuclear burning core is convective. Above the convective radius, the medium is convectively stable and the fluid elements crossing this radius are braked by buoyancy but still have some momentum and they overshoot in the radiative region. The overshoot region undergoes chemical mixing with the underlying H depleted convective layers and this lengthens the MS lifetime of the star. The amount of overshoot (extension of convective motions penetrating into the convectively stable region above) is not known and this is one piece of information which we expect to obtain from the study of such stars as it influences the evolution, the age of such a star, affects the mass-luminosity relation for postMS and evolved stars, etc. In the outer region, because of the large opacity, a very thin convective region exists. There the excitation of the oscillations takes place. For the coolest stars, this is also the region for the main damping process: the interaction turbulent convection–pulsation which can thereby be studied. In-between the energy is transported by radiation.

When the star evolves on the MS, its effective temperature decreases. The outer convective region extends deeper inside the star. On the other hand, as the amount of burning material decreases, the convective core gets smaller and smaller.

Some figures for a typical case can fix ideas: for a $1.7 M_{\odot}$ MS star with 86% central hydrogen left, the inner convective core has a radius $r_c = 0.12 R_* = 136.5$ Mm for a stellar radius $R_* = 1.63 R_{\odot}$. The convective core can be more or less extended depending on the assumed amount of overshoot. The bottom of the upper convective region is found at the radius $r_b = 0.991 R_* = 1124$ Mm.

For an older (*i.e.*, cooler) star (14% central hydrogen left, $R_* = 2.84 R_{\odot}$), the convective core has receded ($r_c = 0.05 R_* = 99$ Mm) while the upper convective region has extended downward ($r_b = 0.915 R_* = 1878$ Mm).

We also encounter δ Scuti stars when they have exhausted their hydrogen and have just moved off the main sequence. Then they have an inert isothermal He core, above which hydrogen burns in a shell.

As a third possibility, stars crossing the instability strip in the δ Scuti domain can be very young stars approaching the Zero Main Sequence (ZAMS). The inner structure of a PMS star crossing the instability strip on its way to the ZAMS, differs from that of an MS star as it does not yet burn hydrogen. The central regions then are radiative. The star lives on its gravitational energy. The outer layers remain very similar to that of an MS δ Scuti star at the same location in the HR diagram, that is, with the same effective temperature and luminosity.

This global and relatively simple picture of the structure of a δ Scuti star is in reality complicated by the fact that these stars are generally fast rotators. Measurements of the surface projected velocities $v \sin i$ show that it can reach values as large as 200 km/s (thus roughly corresponding to a ratio of the centrifugal acceleration to the gravity up to 30–50%) (Royer *et al.* 2002a, b). At these rotational velocities, the centrifugal force distorts quite significantly the star which is no longer spherically symmetric. Theory tells us that rotation generates instabilities which cause both mixing of the chemical elements and transport of angular momentum in the radiative region which in turn affects the rotation profile. How fast is the convective core rotating? How does the rotation–convective interaction affect the structure of the core? Heavy 3D hydrodynamical calculations have recently simulated a rotating convective core for an A-type star with a typical $2 M_{\odot}$ and a rotation equal to the solar rotation which is still rather slow for this type of star (Browning *et al.* 2004). One important result of

interest here is that the convective core is prolate, *i.e.*, elongated towards the poles but the overshoot region keeps a spherical symmetry. The amount of overshoot is found to amount to a few tenths of the pressure scale height which is the expected order of magnitude. More importantly, the overshooting extension depends on the rotation rate and therefore must not be expected to be the same for different types of stars and even from one star to another. It is of course important to have some means of testing such numerical results. The best way, theoretically speaking, is seismology as these stars fortunately oscillate. Hence the idea is to obtain precise seismic information about hydrodynamical processes that take place below the emitting surface, even deep inside the star in order to test the modelling of physical processes which have more general applications and can be applied to other types of stars, namely transport processes: overshoot, time dependent convection, rotation interaction with oscillation and with convection.

2.2 Oscillations

δ Scuti stars are pulsating with periods of a few hours. The mechanism which excites the observed modes is similar to that of a thermal machine or heat engine in that, energy is stored in some peculiar region of the star at compression. Hence the (PdV) work done is positive and this mechanical energy feeds the oscillation. The mode is linearly unstable if this driving overcomes the (radiative) damping. Energy storage at the proper phase is performed through an opacity mechanism (kappa mechanisms, see Shibahashi, these proc.). For δ Scuti stars, the main driving operates in the second helium ionisation region. The kappa mechanism excites low (photometrically detected) as well as intermediate (detected through line profile variations) degrees of l . Radial and nonradial modes with low radial order n have their periods of the same order as the local thermal time scale in the driving region and therefore are efficiently excited. In this low frequency regime, the nature of the excited modes can be complicated: either pure p -modes (pressure fluctuation propagation), or pure g (buoyancy restoring force) or mixed p - and g -modes (modes in avoided crossing or *nearly kissing* modes, see Shibahashi (these proc.)) These latter modes are p -modes in the outer region and g -modes in the inner region; they are then good candidates for probing very inner stellar regions.

However one must be aware that seismology of stars other than the sun cannot be handled the same way. We know much less about these stars: uncertainties in stellar parameters (mass, age, chemical composition) – which are needed to model the structure of the star in the first step of an iterative process – are in most cases too large. Therefore for stars, one must determine all these unknowns together with extracting the information about the physics of the star. This would still be quite easy if we knew for each detected frequency what the associated mode is: and what the degree l , the radial order n and the azimuthal order m are. It is indeed crucial to know which modes we observe as we have to compare with frequencies of the same modes computed from the to-be-probed stellar models. We also need to know the nature of the oscillation mode which we detect and which frequency we measure. In other words, we need to know the resonant cavity which the mode probes as well as its surface geometrical description. This latter one is usually represented by a spherical harmonic of given l , m as the spherical harmonics can define an orthonormal basis and form a very convenient representation of the stellar oscillation (as long as the star is not rotating too fast and

in absence of other processes breaking the axisymmetry). These numbers, n , l , m also have to be determined in the iterative process.

This requires very high quality data, long observation runs, and if possible using different techniques in order to collect as much information as possible – seismic and nonseismic. Far much less ambitious than for the sun, what is now currently being attempted is building a *seismic model* by means of a *best model fitting* method, an optimized procedure which aims at building a model which fulfills all the collected observational constraints including the best fit to the observed oscillation frequencies (in the χ^2 sense). Of course, this model may not be unique but at least it bears some improvement compared to what was known before. In addition, masses, ages for all stars except for the sun, are not known with a high precision. Building seismic models can help to determine these parameters with some accuracy for a large number of stars, in view of statistical studies which are needed for instance, in chemical evolution of galaxies.

3. Seismology of δ Scuti stars: Not an easy task

Seismology of δ Scuti stars, although very promising, has not yet been very fruitful as it faces serious difficulties which will have to be overcome first. If one measures the frequencies, ν_{nlm} , with a high enough precision and if one identifies the associated n , l , m values, even if the stellar parameters are not well known (unlike the sun), the excited modes of δ Scuti stars are discriminant enough that their frequencies enable to determine these stellar parameters. The difficult part, however, is identification of the modes.

3.1 Mode identification

Oscillations of δ Scuti stars are not fully understood: all modes (at least the lowest l degrees) are theoretically expected to be (linearly) excited in a given frequency range but not all modes in this range are detected (see for instance, for the optimal case FG Vir (Breger *et al.* 2005)). Hence the question is: which ones are detected? As an illustration of the diversity of power spectra for δ Scuti and difficulties to interpret them, we can mention 3 stars:

- FG Vir which oscillates with a large number of modes; at least 19 frequencies a few modes seem to have been firmly identified (Breger *et al.* 2004). The star is a moderately fast rotator with $v \sin i = 80$ km/s.

Based on this mode identification, the model fitting method shows that no stellar model can provide frequencies which match simultaneously all the observed frequencies (Breger *et al.* 1999; Daszynska-Daskiewicz *et al.* 2005).

- 1 Mon shows no more than 3 well determined frequencies; two have tentatively been identified (Balona *et al.* 2001).
- HR 6534 (friendly baptised Orianita) illustrates a case of a real rapidly rotating δ Scuti star. In such a case, perturbation theory in order to include the effect of rotation, even at 3rd order, to compute the frequencies can no longer apply (Suarez *et al.* 2005).

Mode identification usually proceeds in determining the degree l of the mode either by using near equidistance in the power spectra (large and small separation), using multicolour photometry (l degree) and line profile variations (m azimuthal order) and

comparing observed and computed frequencies (radial order n). in the case of δ Scuti none of these methods are really easy to use. Indeed, the existence of nearly kissing modes in the observed frequency range, although most interesting for probing the inner structure, destroys the near equidistances as their frequencies do not follow the pure p -mode 'rules'. Modes are also missing and this also perturbs a possible near equidistant distribution. Direct comparison between theoretical and observed frequencies is difficult as several choices of stellar models are possible within the uncertainties in the stellar parameters (mass, effective temperature, luminosity, chemical composition, etc.). Hence the stellar model to be used is not unique.

As the outer layers of δ Scuti stars are convective, identification of the l degree through multicolor photometry has been found very dependent on the adopted convection treatment (for instance on the value of the mixing length in a classical mixing length description) (Balona & Evers 1999) and a reliable time dependent convection description is necessary.

For δ Scuti stars, fast rotation is also a serious problem in the mode identification framework as it not only destroys the equidistant rotational splitting between $m \neq 0$ modes but for the fastest cases, the description with one single spherical harmonic is no longer valid, hence attempting to identify one mode with one single (n, l, m) set of numbers no longer makes real sense.

3.2 Linear excitation and damping: Time dependent convection

The linear time dependence of the oscillation in a given mode, either displacement or temperature changes for instance, is expressed as $ae^{i(\omega t + \kappa t)}$ where ω is the pulsation frequency, κ the linear growth rate and a is a complex number representing the complex amplitude of the oscillation which is underdetermined (or arbitrary) in a linear description.

While we understand the excitation mechanism for δ Scuti stars (kappa mechanism), and are able to predict roughly the blue side of the instability strip ($\kappa > 0$), the return to stability on the red edge ($\kappa < 0$) is attributed to an efficient interaction between convection and pulsation and has been quantified with some confidence only recently (see section 4.1).

3.3 Finite amplitude oscillations

When a mode is linearly unstable (excited), its amplitude $|a|$ can grow with time (roughly on a thermal time scale) from an infinitesimal perturbation to a finite value which, when larger than the detection threshold value, is the observed value. It would help the mode identification if one was able to estimate theoretically the amplitude of each mode and to determine whether each mode is detectable or not. Unfortunately, the saturation mechanism which stops the mode amplitude increase is not definitely identified in the case of δ Scuti stars.

The amplitude growth can be limited by a self saturation mechanism (energy of the mode is pumping back to the stellar thermal reservoir of the star) or by nonlinear resonant coupling with stable modes (energy transfer from unstable modes to stable ones). The self saturation mechanism seems to lead to too large amplitudes compared with the observations. Several types of resonant coupling are possible. The most probable resonant cases involve two mode or three resonant mode couplings but for a given

unstable mode, many stable modes can be resonantly coupled to a given unstable mode and the estimation of its amplitude is of a statistical nature, *i.e.*, hardly predictable for individual modes (Dziembowski *et al.* 1988).

To proceed further in the understanding and modelling of the finite amplitude behavior, we need some observational hints and we must await the detection – from future (ground based and space) observations – of a much larger number of modes for each individual star as in the case of FG Vir (Breger *et al.* 2005).

4. Advances in theoretical modelling

4.1 Time dependent convection

As already mentioned earlier, the basic mechanism which makes the modes linearly unstable is the kappa mechanism. On the red side of the instability strip however, the growth rates are not properly computed when the convection is assumed frozen in. Indeed, the frozen-in approximation, which assumes that the convective flux does not vary over a pulsation period ($\delta F_c = 0$), is no longer valid because the eddy turn-over time scale is no longer much larger than the pulsation period. Frozen convection models do not predict the stabilization of the modes (growth rates κ negative) at the red edge of the instability strip. This stabilization (required because a red edge is observed) can only be obtained by time dependent convection. Modelling over the whole extension of the instability strip thus requires to include a time dependent convection (TDC) in the theoretical description.

Several approaches have been developed in order to provide a 1D description of 3D time dependent convection and its interaction with pulsation. They differ in their assumptions and simplifications (Houdek *et al.* 1999; Houdek 2000; Xiong *et al.* 1997; Xiong & Deng 2001; Dupret *et al.* 2004). As a consequence, the temperature for the red edge of the fundamental radial mode for instance is not quite exactly the same depending on the adopted description, one obtains for instance T_{RE} (Xiong & Deng 2001) $< T_{RE}$ (Dupret *et al.* 2004) $< T_{RE}$ (Houdek 2000).

Nevertheless, these approaches agree on some issues such as the following:

- Main driving and damping always occur around the second ionization zone of H_e , because the periods are of the same order as the thermal time-scale in this region. Therefore, the bottom of the convective region in the outer layers must reach down the H_e^{++} region for damping by convection to be efficient in δ Scuti stars ($|\delta F_c| > |\delta F_{rad}|$). This implies a strong dependence of the red edge location on the effective temperature and on the mixing-length parameter α (related to the convective zone size).
- On the cool side of the instability strip and at the bottom of the convective envelope δF_{rad} drops outward rapidly. In the frozen convection description, this implies that energy is blocked at the hot phase of pulsation, which produces a positive PdV work (see in Fig. 1 the increase of the work for the FC case near the convection zone bottom). In the TDC case, δF_c is sufficient to produce a negative work that counterbalances this effect, leading to the stabilization of the modes (see in Fig. 1 the work due to δF_c (W_{Fc})).
- Several mechanisms contribute to the driving and damping. The work integral can be split into $W_{tot} = W_{FR} + W_{Fc} + W_{pt} + W_{Ect} + W_{visc}$ where W_{FR} is the contribution to the work due to the radiative flux variations, W_{Fc} comes from the convective

flux variations, W_{pt} comes from the turbulent pressure variations, W_{Ect} comes from turbulent kinetic energy dissipation variations and W_{visc} from turbulent viscous stress variations.

Hence the red edge is strongly dependent on the mixing length α , on the metallicity, Z , and the model atmosphere used in the computation. In turn, these quantities can be thereby determined or constrained by the observations.

Using the description of Dupret *et al.* (2004) for instance, a stellar model with a $1.8\text{--}2\text{ M}_{\odot}$ mass with a temperature of about 6500 K sees its modes between g_3 up to p_7 – and beyond – unstable in the frozen convection approximation whereas in the TDC approximation all these modes are stable and such a star is no longer a pulsating star. Figure 1 shows the impact of including the interaction between convection and oscillation on the sign of the work integral.

A consensus is yet to be obtained. For instance Houdek (2000) finds that the perturbation of the turbulent pressure δ_{pt} is the main stabilizing agent; Xu *et al.* (2002) find instead that δ_{pt} rather contributes to excitation; Dupret *et al.* (2004, 2005a) find that W_{Ect} balances W_{pt} hence $W_{tot} \simeq W_{Fc} + W_{FR}$.

4.2 Mode identification

We only consider here photometric methods for the identification of the degree l (see also Handler, these proc.). Photometric variations in a non-radial pulsator depend on various contributions. One comes from the effect of effective temperature variations on the flux variations; a gravity term describes the effective gravity variations (including the acceleration of the comoving frame) on the flux variations; and finally a geometrical term measuring the variation of the visible surface of the star. This last term is very important, as it is proportional to $(1 - l)(l + 2)$. It thus allows the identification

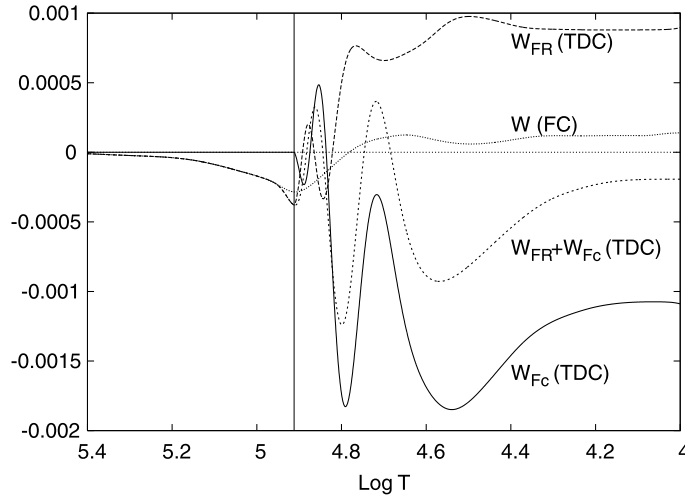


Figure 1. Different physical components of the work integral as a function of $\log T$, obtained with FC and TDC treatment for the radial p_3 mode of a model with $1.8\text{ M}_{\odot}$ and $T_{\text{eff}} = 6680\text{ K}$. W_{FR} (resp. W_{Fc}) is the contribution to the work due to the radiative flux (resp. convective flux variations). The vertical line is the base of the convective envelope. On the instability strip red side, the time variation of the convective flux plays a significant role in all the convective envelopes.

of l . The usual method for this identification is based on observations in multicolour photometry. We can write: $m_\lambda = A_\lambda \cos(\omega t + \phi_\lambda)$ where λ represents the wavelength or more practically the observed wavelength passband. The weights of the different physical terms (see above) depend on the wavelength passbands. Therefore, comparison between the theoretical and observed amplitude ratios $A_{\lambda_1}/A_{\lambda_2}$ and phase differences $\phi_{\lambda_2} - \phi_{\lambda_1}$ makes the identification of l possible (for a review see Garrido 2000). However, a careful, physical and numerical, description of the atmosphere (monochromatic flux and limb-darkening derivatives with respect to temperature and gravity) is required (Garrido 2000; Dupret *et al.* 2003; Barban *et al.* 2003). Also, the theoretical predictions are very sensitive to the amplitude and phase of effective temperature variations, hence a proper description of the oscillation energetics in the superficial layers is required (Dupret *et al.* 2003). For δ Scuti stars, time-dependent convection models are necessary (Dupret *et al.* 2005b). In turn, this enables to probe inefficient convection treatments in subphotospheric layers as oscillating quantities which are strongly dependent on convection can be directly determined from observations (Moya *et al.* 2004; Daszynska-Daskiewicz *et al.* 2005).

Rotation can also perturb on a significant level the mode identification as was pointed out by Soufi *et al.* (1998) and shown quantitatively by Daszynska-Daskiewicz *et al.* (2002).

4.3 Rotation

Fast rotation by distorting the structure of the star affects its oscillations in various ways. These effects are no longer negligible whenever the rotational velocity v reaches ~ 100 km/s–200 km/s for the excited δ Scuti-like oscillation modes. Interaction between rotation and pulsation has therefore been included to compute the oscillation frequencies corrected for rotation effects (Dziembowski & Goode 1992; Soufi *et al.* 1998; Goupil *et al.* 2000; Suarez *et al.* 2005; Karami *et al.* 2005) and linear growth rates (Lee & Baraffe 1995) using perturbation theory and/or simplifying approximations. These developments enable quantitative studies and are included in best model fitting methods. For the fastest stars, however, one will have to turn to nonperturbative treatments of the pulsation–rotation interaction which are being currently developed (Rieutord & Valderatto 1997; Lignières *et al.* 2001).

5. Pre-main sequence δ Scuti stars

Pre-main sequence A–F pulsating stars should be precursors of δ Scuti stars or roAp stars. It is of great interest to determine the nature of their oscillations and their internal properties, particularly their rotation profiles as we could then obtain constraints about the evolution of the internal rotation from well before arrival on the main sequence.

Because of their relatively simple inner structure, no nearly kissing modes are expected in the frequency range where observed modes are detected. Hence one expects simple patterns (near equidistance) in the power spectra of these stars at least for the $m = 0$ modes. However, their usually fast rotation certainly complicates the structure of the frequency distribution in their power spectra.

Observationally, although the first claim of the observation of a variable PMS δ Scuti star goes back to 1972 (Breger 1972), it was not until 1995 after the detection of the variability of HR 5999 (Kurtz & Marang 1995) that the field really developed.

More than a dozen such stars are now known (see for a review Catala 2003; Marconi & Palla 2004).

We illustrate this topic with two interesting cases:

5.1 An eclipsing double line spectroscopic binary PMS system: RS Cha

The eclipsing double line spectroscopic binary system RS Cha is known as a PMS system since 2000 (Mamajek *et al.* 2000). As the orbit is known, masses, radii for both components are known with high precision. The two stars are of similar masses slightly larger than $1.8 M_{\odot}$. Inclination angle and $v \sin i$ are well determined. In addition, photometric determinations provide the effective temperatures (T_{eff}) and luminosities (L) of both stars (A, B). Spectroscopic data at SAAO have been analysed and yield the metallicity $Z = 0.028 \pm 0.004$ (Alecian *et al.* 2005).

With the knowledge of these stellar parameters, it is found that it is possible to model simultaneously both stars at same age. Hence to obtain more stringent constraints on the physical description of this type of stars, a better determination of T_{eff} is necessary on one side and more information about the pulsations of both stars. From residual orbital radial velocities obtained from the SAAO 2004 data, it is definitely confirmed that both stars oscillate (Alecian *et al.* 2005). The data quality unfortunately is not good enough to provide secure periods yet.

5.2 A binary PMS system HD 104237

We now turn to another binary PMS star system: HD 104237 (friendly baptised Christianita). The A component appears to be a δ Scuti type pulsator showing multiperiodic pulsations (5 frequencies + 3 possible) (Donati *et al.* 1997; Kurtz & Müller 1999; Böhm *et al.* 2004).

From the orbit, the mass ratio is known with some precision. A photometric determination for the primary gives $T_{\text{eff}} \sim 8500$ K; $\log L/L_{\odot} \sim 1.55$ although other determinations show that these values are not secure. Nevertheless using these values for the purpose of illustration, Fig. 2 shows a comparison between the observed frequencies in a power spectrum and the computed frequencies for a stellar model satisfying these constraints (see Samadi *et al.*, these proc.). Computed frequencies for all modes with $l = 0$ to $l = 3$ frequencies in the same frequency range as in observations, are included.

When no rotation effect is included, the computed power spectrum displays a simpler pattern than for classical (MS, off MS) δ Scuti stars. The excited modes for that particular case are of rather high radial order, definitely non-radial modes. They display a near equidistant pattern which should provide a determination of the large separation. Not all these computed modes are detected however. Besides the pattern is more complicated when one takes into account the m-splitting due to rotation. Here we have taken a typical rotation $v = 50$ km/s compatible with the observed $v \sin i = 12 \pm 2$ km/s.

This example shows that one faces the same problems as for MS δ Scuti stars although less severely. To go further, one nevertheless needs mode identification; a determination of the rotation period would definitely help. The next step is a stability analysis in order to provide additional constraints on the modelling of this PMS δ Scuti like star.

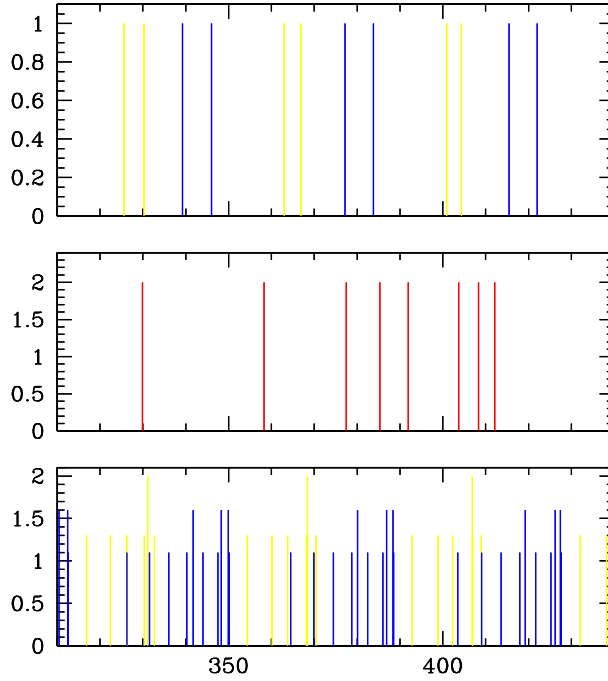


Figure 2. Top: Computed power spectrum with $l = 0 - 3$, $m = 0$ mode frequencies (arbitrary amplitudes); no rotation effect included. Middle: Observed frequencies for HD 104237 (arbitrary amplitudes). Bottom: same as top but with $m \neq 0$ modes included (represented by smaller amplitudes) with a chosen rotation velocity $v = 50$ km/s.

Acknowledgements

RS and MJG acknowledge financial supports by Programme National de Physique Stellaire (PNPS, France) and by the Aryabhata Research Institute of Observational Sciences (ARIES, Nainital, India). We thank C. Barban and F. Baudin for fruitful discussions. We thank the referee G. Handler for his relevant suggestions.

References

- Alecian, E., Catala, C., Van't Veer, C. *et al.* 2005 (in press).
 Balona, L. A. & Evers, E. A. 1999, *MNRAS*, **302**, 349.
 Balona, L. A., Bartlett, B., Caldwell, J. A. R. *et al.* 2001, *MNRAS*, **321**, 2391.
 Barban, C., Van't Veer, C., Goupil, M. J. *et al.* 2003, *A & A*, **405**, 1095.
 Böhm, T., Catala, C., Balona, L., Carter, B. 2004, *A & A*, **427**, 907.
 Breger, M. 1972, *ApJ*, **171**, 539.
 Breger, M., Rodler, F., Pretorius, M. L. *et al.* 2004, *A & A*, **419**, 695.
 Breger, M., Pamyatnykh, A. A., Pikall, H., Garrido, R. 1999, *A & A*, **341**, 151.
 Breger, M. 2005, In 'The A-Star puzzle', IAU Symp. 224, Cambridge Univ. press, p. 335.
 Breger, M., Lenz, P., Antoci, V. *et al.* 2005, *A & A*, **435**, 955.
 Browning, M. K., Brun, A. S., Toomre, J. 2004, *ApJ*, **601**, 512.
 Catala, C. 2003, *ApSS*, **284**, 53.
 Daszynska-Daskiewicz, J., Dziembowski, W. A., Pamyatnykh, A. A., Goupil, M. J. 2002, *A & A*, **392**, 151.
 Daszynska-Daskiewicz, J., Dziembowski, W. A., Pamyatnykh, A. A. 2005, *A & A*, (in press).

- Donati, J. F., Semel, M., Carter, B. D. *et al.* 1997, *MNRAS*, **291**, 658.
- Dupret, M.-A., De Ridder, J., De Cat, P. *et al.* 2003, *A & A*, **398**, 677.
- Dupret, M.-A., Grigahcène, A., Garrido, R. *et al.* 2004, *A & A*, **414**, L17.
- Dupret, M.-A., Grigahcène, A., Garrido, R. *et al.* 2005a, *A & A*, **435**, 927.
- Dupret, M.-A., Grigahcène, A., Garrido, R. *et al.* 2005b, *MNRAS*, **361**, 476.
- Dziembowski, W. A. 1990 In *Progress of Seismology of the Sun and Stars. Lecture Notes in Physics*, Springer-Verlag, **367**, 359.
- Dziembowski, W. A., Goode, P. R. 1992, *ApJ*, **394**, 670.
- Dziembowski, W. A., Krolikowska, M., Kosovitch, A. 1988, *Acta Astron.*, **38**, 61.
- Garrido, R. 2000, In: *The 6th Vienna Workshop on δ Scuti and related stars, PASP Conf. Ser.*, **210**, 67.
- Goupil, M. J., Dziembowski, W. A., Pamyatnykh, A. A., Talon, S. 2000, *ASP Conf. Ser.*, **210**, 267.
- Goupil, M. J., Talon, S. 2002, *ASP Conf. Ser.*, **259**, 306.
- Houdek, G., Balmforth, N. J., Christensen-Dalsgaard, J., Gough, D. 1999, In “*Stellar Structure: Theory and test of convective energy transport*”, *ASP Conf. Ser.* **173**, 317.
- Houdek, G. 2000, In: *The 6th Vienna Workshop on δ Scuti and related stars, PASP Conf. Ser.*, **210**, 454.
- Karami, K., Christensen-Dalsgaard, J., Pijpers, F. P. *et al.* 2005, *A & A*, (in press).
- Kurtz, D. W. 2003, *Astrophysics and Space Science*, **284**, 29.
- Kurtz, D. W., Marang, F. 1995, *MNRAS*, **276**, 191.
- Kurtz, D. W., Muller, M. 1999, *MNRAS*, **310**, 1071.
- Lee, U., Baraffe, I. 1995, *A & A*, **301**, 419.
- Lignières, F., Rieutord, M., Valdettaro, L. 2001, SF2A, EdP-Sciences, Conf. Ser. in Astronomy and Astrophysics, p. 127.
- Mamajek, E. E., Lawson, W. A., Feigelson, E. D. 2000, *ApJ*, **544**, 356.
- Marconi, M., Palla, F. 2004, In “*The A-star puzzle*”, IAU 224, Cambridge Univ. Press, p. 69.
- Moya, A., Garrido, R., Dupret, M. A. 2004, *A & A*, **414**, 1081.
- Pamyatnykh, A. A. 2003, *Astrophysics and Space Science*, **284**, 97.
- Poretti, E., Garrido, R., Amado, P. J. *et al.* 2003, *A & A*, **406**, 203.
- Rieutord, M., Valdettaro, L., 1997, *J. Fluid Mech.*, **341**, 77.
- Royer, F., Gerbaldi, M., Faraggiana, R., Gomez, A. E. 2002a, *A & A*, **381**, 105.
- Royer, F., Grenier, S., Baylac, M. O. *et al.* 2002b, *A & A*, **393**, 897.
- Suarez, J. C., Goupil, M. J., Morel, P. 2005, *A & A*, (in press).
- Soufi, F., Goupil, M. J., Dziembowski, W. A. 1998, *ApJ*, **334**, 911.
- Xiong, D. R., Cheng, Q. L., Deng, L. 1997, *ApJS*, **108**, 529.
- Xiong, D. R., Deng, L. 2001, *MNRAS*, **324**, 243.
- Xu, Y., Li, Z. P., Deng, L. C., Xing, D. R. 2002, *Chinese J. Astron. Astrophys.*, **2**, 441–448.

Pulsations in Subdwarf B Stars

C. Simon Jeffery

Armagh Observatory, College Hill, Armagh BT61 9DG, Northern Ireland.

Abstract. Subdwarf B stars play a significant role in close binary evolution and in the hot star content of old stellar populations, in particular in giant elliptical galaxies. While the question of their origin poses several problems for stellar evolution theory, one of their most fascinating properties is the presence of multi-periodic 2–3 minute oscillations. Interpreting these oscillations optimally requires the correct identification of the modes. Partial identifications can be obtained using high-speed observations of radial velocity and colour variations. We review some of the several attempts to make such observations, most recently with the Multi-Site Spectroscopic Telescope campaign and with ULTRACAM.

Key words. Stars: oscillations—stars: variables: other (EC14026/sdBV)—stars: individual: KPD2109+4401, PB8783, PG1605+072, HS0039+4302, PG0014+067.

1. Introduction

An examination of a false-colour ultraviolet photograph of the globular cluster NGC2808 reveals a number of blue stars too bright to be white dwarfs, and too faint to be upper main-sequence stars (Brown *et al.* 2001). Similarly, the ultraviolet spectra of giant elliptical galaxies, although devoid of massive hot stars, reveal a flux excess that cannot be explained by the visible cool-star population (Brown *et al.* 2000). It is clear that old stellar systems are dominated by a population of subluminous hot stars.

The subluminous B (or sdB) stars were first identified as an important spectral class by Greenstein & Sargent (1974). In their survey of faint blue stars in the Galactic halo, they demonstrated that, on the Hertzsprung–Russell diagram, the sdB stars lie between the upper main sequence and the white dwarf sequence, possibly on a blue extension of the horizontal branch similar to that observed in globular clusters. This was confirmed by subsequent fine analyses of sdB stars in the field and extremely blue horizontal-branch stars in globular clusters (Heber *et al.* 1984; Heber 1986; Heber *et al.* 1986). These studies established that sdB stars are core-helium burning stars of $\sim 0.5 M_{\odot}$ with a very thin surface layer of hydrogen. With effective temperatures $25\,000 \leq T_{\text{eff}}/\text{K} \lesssim 35\,000$, the atmospheres of these stars are entirely radiative. Combined with high surface gravities ($5 \lesssim \log g \lesssim 6$), the radiative atmospheres provide ideal conditions for the competition between radiative levitation and gravitational settling to act differentially on different atoms. While levitation is governed by the radiative cross-section of the dominant ion of a given atomic species, the atomic mass

governs the gravitational force. The resulting diffusion affects the chemical stratification in the atmosphere. In sdB stars, the diffusion timescale ($\sim 10^5$ y) is short compared with evolution timescales, thus leading many sdB stars to appear depleted in helium (Michaud *et al.* 1989) and to demonstrate other unusual abundance characteristics (O'Toole 2004).

The sdB stars represent the quintessential horizontal branch star problem: “How can a star lose virtually all of its hydrogen-rich outer layers before it ignites helium reactions in its core?” Originally thought to be mainly single stars, an increasingly large fraction of sdBs are now known to be binaries. The companions are either cool main-sequence stars (e.g., Allard *et al.* 1994; Aznar Cuadrado & Jeffery 2001) or white dwarfs (Maxted *et al.* 2000). This provides mass transfer by Roche lobe overflow near to the tip of the first giant branch, *i.e.*, just before core-helium ignition, as a natural mechanism for removing the outer hydrogen. It remains important to verify this hypothesis by measuring their core and envelope masses directly, to explain in detail the general properties of the sdB stars and to account for exceptional cases. Indeed, there is evidence that up to a third of sdBs may be single. It has been suggested that these were formed as the result of a merger of two helium white dwarfs (Iben & Tutukov 1985, 1986; Iben 1990; Saio & Jeffery 2000).

2. Discovery of sdB pulsations

The discovery of pulsations in sdB stars was almost fortuitous. Kilkenney *et al.* (1997) had been searching for variability in a number of white dwarf candidates including one EC 14026–2647. Finding short-period low-amplitude variability they investigated further and discovered that the star was an sdB star with an F-type companion. Extended photometric monitoring demonstrated a light curve with at least two independent frequencies which could only be attributed to a non-radial oscillation in one or the other but most likely the sdB star. Subsequently, photometric monitoring of a much larger sample of sdB stars led to the discovery of several more EC14026–2647 or sdB variables (sdBVs), although the non-variables far outnumbered the variables. The second to be analysed was PB 8783 (Koen *et al.* 1997). Following the discovery paper, an extended multi-site photometric campaign was able to resolve a spectrum of closely-spaced frequencies between 7 and 10 mHz (O'Donoghue *et al.* 1998), indicating that several non-radial oscillations were being excited at the same time.

Meanwhile, and independently of the SAAO group, Charpinet *et al.* (1996, 1997) had been making theoretical calculations which predicted that some sdB stars should pulsate. They had been searching for sdB pulsators with little success until they observed PG1047+003 (Billères *et al.* 1997) and KPD 2109+4401 (Billères *et al.* 1998). Meanwhile the latter star had formed part of the SAAO survey; Koen *et al.* (1998) published its power spectrum in the same year. Both groups obtained nearly identical light curves, with a clutch of frequencies between 5 and 6 mHz.

Many sdBVs were discovered from photometry covering only one or two hours. This is an effective way of establishing whether oscillations with periods of 100 to 500 seconds (2–10 mHz) are present, but selects against longer period oscillations. More careful observations covering many hours have recently established the presence of a second-group of pulsating sdB stars with periods of 1800s and upwards (< 0.5 mHz; Green *et al.* 2003). These long-period sdBVs are almost universally cooler than the short-period sdBVs.

The increasing number of sdBVs now known, their frequency spectra and their general properties are reviewed elsewhere (O'Donoghue 1999; Charpinet *et al.* 2001; Kilkenny 2002).

3. Excitation mechanism

Stellar pulsations are found in stars distributed widely across the Hertzsprung–Russell diagram (Gautschi & Saio 1995, 1996). Radial pulsations are driven by instabilities related to opacity arising from the ionization of different ions. Thus δ Cepheid and RR Lyrae pulsations, *inter alia*, are driven by He II ionization, whilst β Cepheid pulsations are associated with ionization of iron-group elements (Fe-bump opacities). Even hotter stars – GW Vir variables – are associated with highly ionized species of carbon and oxygen (CO-bump) (Saio 1996). Each of these driving zones gives rise to an instability strip.

Still other mechanisms produce pulsating variables elsewhere in the HR diagram. Examples include roAp, δ Scuti, γ Doradus, 53 Persei and Luminous Blue variables. Most of these stars pulsate in non-radial modes of low radial order (k) and low spherical degree (l). The sdBVs represent one of these groups. They may lie roughly on the low-luminosity extension of the Fe-bump instability strip. Fe-bump excitation, possibly in a region just beneath the stellar atmosphere enriched in metals as a consequence of chemical diffusion, is the most promising interpretation for their origin (Charpinet *et al.* 1997). However, since both non-radially oscillating and non-variable sdB stars occur with very similar overall dimensions (T_{eff} , $\log g$), the extent and mechanism by which the atmosphere becomes stratified needs to be understood more clearly.

The importance of studying non-radial oscillations lies in their ability to provide information about the stellar interior – this is the science of asteroseismology. For example, oscillations in hot white dwarfs can be used to study the stellar mass (period spacing), the depth and composition of the outer layers (mode trapping), luminosity (both of the above), rotation speed and magnetic field (mode frequency splitting), and evolution time scales (period changes) (Vauclair 1996). To make the most of the observations, it is helpful to be able to identify the mode of each oscillation. These modes are generally spherical harmonics denoted by three integers, k – radial order, l – spherical degree and m – azimuthal wavenumber. Due to cancellation, it is usually only possible to observe low-degree modes ($l, m \lesssim 4$) in stars other than the sun, and it is normally difficult to disentangle these from the light curve alone. Mode identification generally requires information from other sources, e.g., colours, velocities and line profiles. This has motivated the work described in subsequent sections.

4. Radial velocity observations

Immediately after the discovery of sdBV pulsations, we recognized the potential for spectroscopy to contribute to the subject. Mode identification from single-band photometry alone is challenging, and could be assisted by the identification of line-profile variations, whilst the comparison of radial and light amplitudes could be used to determine stellar radii directly. Spectroscopy might also demonstrate the presence of higher-order modes not detected photometrically. The limitations are that the periods are short (100–500s) compared to conventional CCD readout times, the stars are

faint (> 11 th mag.), and the photometric amplitudes are typically only a few tenths of one per cent. High-resolution high-S/N multi-line studies such as those obtained for non-radial oscillations in rapidly rotating bright O and B stars (e.g., Reid *et al.* 1993; Telting *et al.* 1997) would not appear to be feasible. However, the development of new techniques offered the possibility to acquire high-speed spectroscopy of sdBVs.

4.1 1998: PB8783 and KPD2109+4401

Observations of two bright sdBVs PB8783 and KPD2109+4401 were obtained in 1998 October with the 4.2 m William Herschel Telescope using the blue arm of the intermediate dispersion spectrograph ISIS. The R1200B grating and the TEK1 CCD with 1124^2 $24\ \mu\text{m}$ pixels yielded an instrumental resolution (2 pixel) $R = 5000$ in the wavelength interval 4020–4420 Å. This wavelength region was chosen because it contains two strong Balmer lines and, potentially, a number of neutral helium and minor species lines that are normally observed in early-B stars. It also maximises the photon collection rate. The CCD was read out in low-smear drift mode (Rutten *et al.* 1997), in which only a small number j of CCD rows (parallel to the dispersion direction) are read out at one time. A dekker is used to limit the slit-length, thus only a fraction of the CCD window, $\simeq j$ rows, is exposed at one time. After exposing for a short interval, typically 10s, the CCD contents are stepped down by j rows. Each set of j rows is accumulated into a data cube containing n individual 2D spectra, stacked adjacent to one another. More complete details of the instrumental setup and data reduction have been given by Jeffery & Pollacco (2000). The significance of these data was that they provided high time resolution and good wavelength stability with reasonable photon statistics. Conventional methods would have, at that time, resulted in unacceptable dead time due to the read-out time for the CCD. The final data product is a set of files

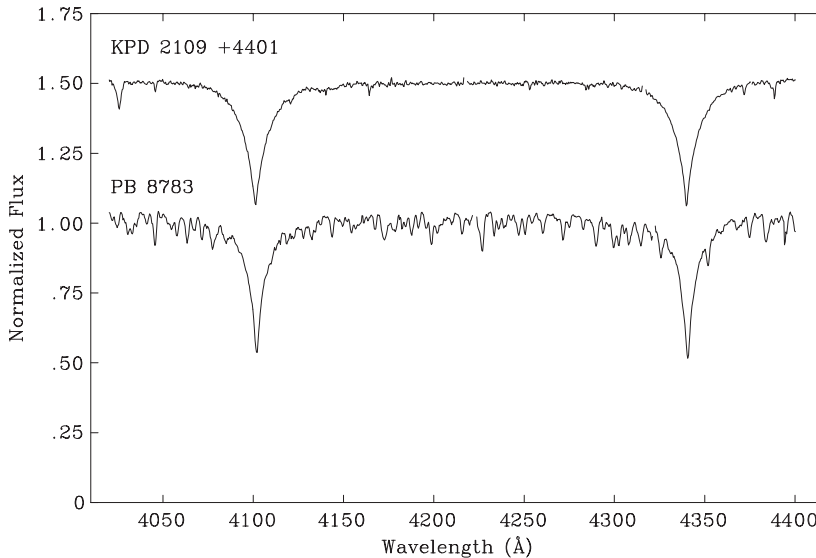


Figure 1. Template spectra of sdBVs KPD 2109+4401 and PB 8783 formed by coadding several hundred short-exposure spectra (Jeffery & Pollacco 2000). Note the broad Balmer lines due to the sdB star in both cases, and the sharp metallic line spectrum of the companion F star in PB 8783.

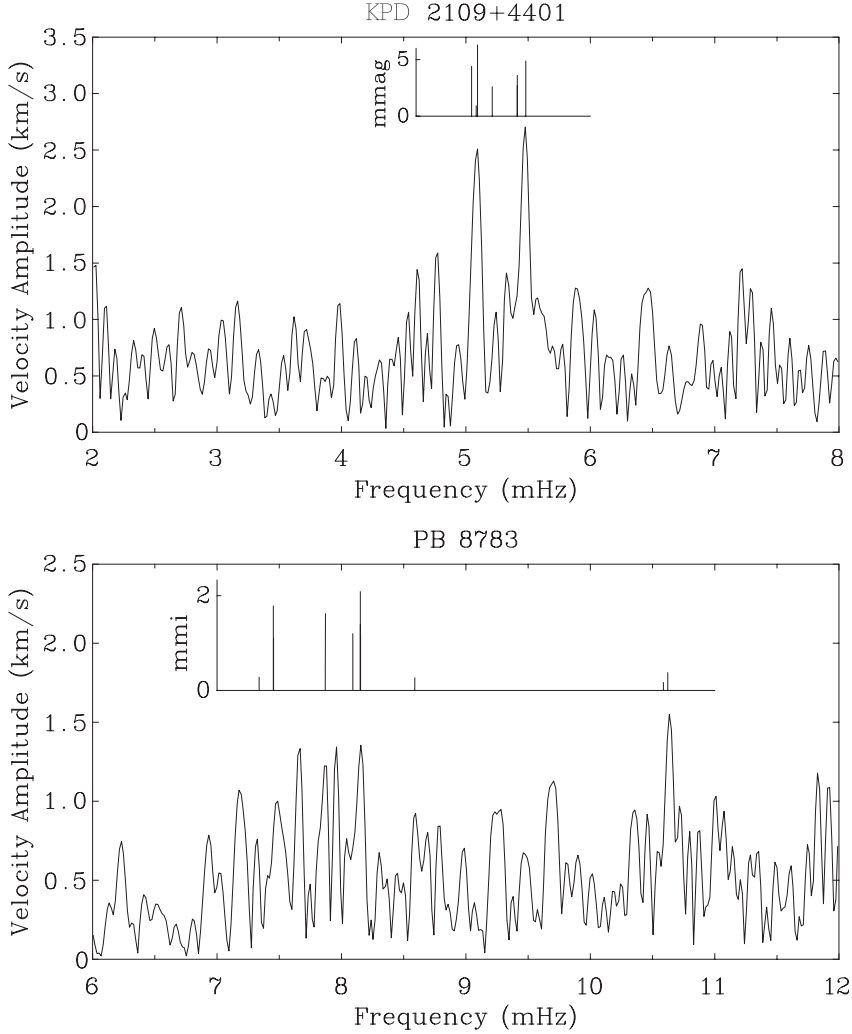


Figure 2. Amplitude spectra from WHT radial velocity observations and schematic representations of the photometry amplitude spectra (inset) for sdBVs KPD2109+4401 and PB8783 (from Jeffery & Pollacco 2000).

each containing n 1-dimensional wavelength-calibrated flux-normalized spectra. Each spectrum is time-tagged.

A template spectrum was constructed by coadding all of the data in a time series (Fig. 1). Individual velocities $v(t)$ were measured by cross-correlation with the template. The functions $v(t)$ were then investigated for periodic content by means of their amplitude spectrum (or discrete Fourier transform) $\mathcal{F}$.

The amplitude spectra $\mathcal{F}$ are shown in Fig. 2, together with a representation of the principal frequencies and amplitudes¹ discovered photometrically $\mathcal{P}$ (O'Donoghue

¹The term amplitude applied here to periodic signals refers to the semi-amplitude a of the sine function $a \sin(2\pi \nu t + \phi)$.

et al. 1998; Koen 1998). In this particular experiment, the information required from the amplitude spectrum is not the significance of the peaks but rather the amplitude associated with expected frequencies. Over the entire range $0.1 \sim 2/\Sigma n\delta t < \nu/\text{mHz} < 1/\delta t \sim 90$, $\mathcal{F}$ shows many peaks of comparable amplitude which on their own have little statistical significance. However the coincidence in Fig. 2 of the highest peaks in $\mathcal{F}$ with the highest peaks in $\mathcal{P}$ is more remarkable. The resolution $2/n\delta t \sim 0.1 \text{ mHz}$ in $\mathcal{F}$ is insufficient to resolve the fine structure observed in much longer photometric time series.

On closer inspection we found that the cross-correlation function for PB 8783 consisted of a broad and a narrow component. The narrow component, corresponding to the metal lines arising in the cool companion, had a mean velocity trend ($b(t)$) opposite to that of the broad component, corresponding to the Balmer lines in the sdB star. There were no high frequency components in the cool star ccf. While the low-frequency drifts $b(t)$ could be due to mutual acceleration in the binary system, with a period of roughly 0.5–3 days, the oscillations are clearly seen to arise in the sdB star and not in the F star.

4.2 2000: PG1605+072

O'Toole *et al.* (2000) had announced preliminary results of independent radial velocity observations of the brightest sdBV PG1605+072. This star has both the largest number of observed frequencies (> 50 , Kilkenney *et al.* 1999) and the largest amplitudes ($> 0.025 \text{ mag}$), making it an ideal target for asteroseismological study. O'Toole *et al.* found a 14 km s^{-1} amplitude periodic variation at the principal frequency of 2.10 mHz from a low-dispersion campaign carried out over 14 nights at a 2 m telescope with a duty cycle for individual observations of $\sim 60\text{--}75\text{s}$.

Therefore during 2000, we observed PG1605+072 in the same way as before using the WHT and the Anglo-Australian Telescope (Stathakis & Johnston 1997) in an effort to obtain near continuous spectroscopy with a duty cycle of $\sim 13\text{--}22\text{s}$. Regrettably, the original plan to observe this target for nearly 72 h was reduced to something like 16 h. Datasets from both telescopes were combined and analysed as before (Woolf *et al.* 2002). A section of the power spectrum is shown in Fig. 3, with a schematic representation of the photometric power spectrum superimposed (Kilkenney *et al.* 1999). While amplitudes of weaker modes could now be measured, the frequency resolution of the dataset was limited.

4.3 2002: PG1605+072: MSST

Following these campaigns it was clear that progress in understanding the radial velocity amplitudes of the principal modes in sdBVs would only be obtained by substantially longer observing runs. Thus, in May 2002, a large consortium set out to obtain 4 m high-speed spectroscopy of PG1605+072 with 24 h coverage over several nights. This campaign was strongly supported by spectroscopy from 2 m telescopes, and by photometry from a large network of smaller telescopes, including the Whole Earth Telescope. The 2 m spectroscopy and the photometry campaigns were extremely successful; preliminary results have been reported by Heber *et al.* (2003) and O'Toole *et al.* (2004). Regrettably, the 4 m campaign was again severely limited by weather and telescope allocations and instrumental problems. Formal reports will be published shortly.

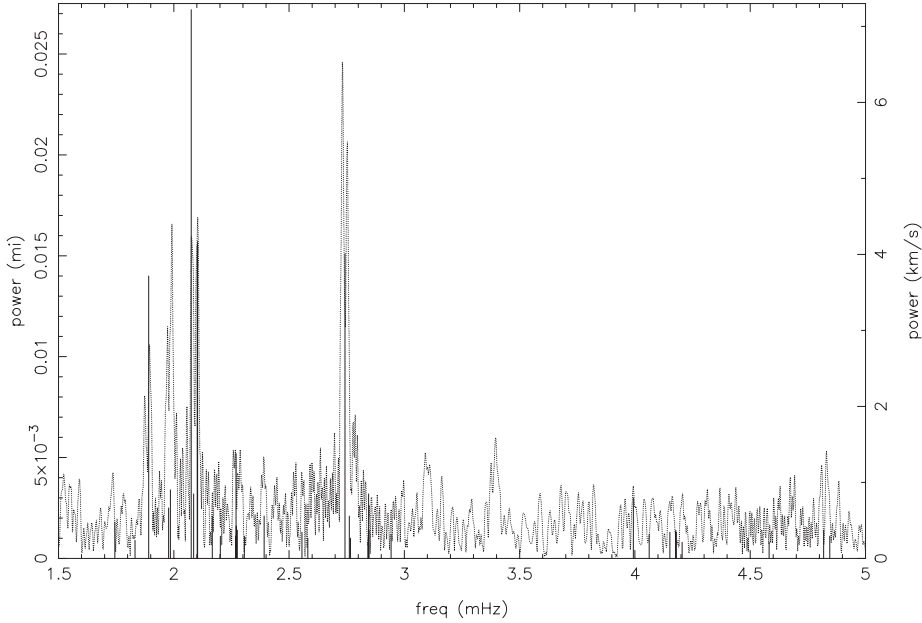


Figure 3. Amplitude spectra from WHT and AAT radial velocity observations (dotted line, labelled right) and schematic representation of the photometry amplitude spectra (solid line, labelled left) for sdBV PG1605+072 (from Woolf *et al.* 2002).

5. Multicolour high-speed photometry

An alternative and much more powerful approach to the partial identification of non-radial modes is the amplitude ratio method (Heynderickx *et al.* 1994). This makes use of the fact that the ratio of the photometric amplitude at one wavelength to the amplitude at another wavelength (a_λ/a_{λ_0}) is a function of the spherical degree l of the oscillation and is independent of the inclination of the pulsation axis i and the azimuthal wavenumber m . The method is simplified for use in broadband multicolour photometry by using the ratio of the amplitudes in different wavebands (e.g., $a_{r'}/a_{u'}$ if using Sloan Digital Sky Survey filters). Koen (1998) had already made an attempt to identify modes in KPD2109+4401. While indicative, his results showed room for improvement in terms of time-resolution and photometric precision.

ULTRACAM is a high-speed 3-channel CCD camera. Dichroics split the light up into ultraviolet, visual and red wavelengths which are simultaneously imaged with frame transfer CCDs (1024 by 1024 imaging area). These CCDs use a masked area to shunt exposures rapidly from the imaging area, which can then be read out while the next exposure is running. Mounted on the 4.2 m William Herschel Telescope, the combination of high throughput and high time resolution provides an outstanding new tool with which to explore pulsating sdB stars.

5.1 2002: KPD2109+4401 and HS0039+4302

The first observations were obtained in 2002 September for the targets KPD2109+4401 and HS0039+4302. A portion of the light curves for each of these targets is shown

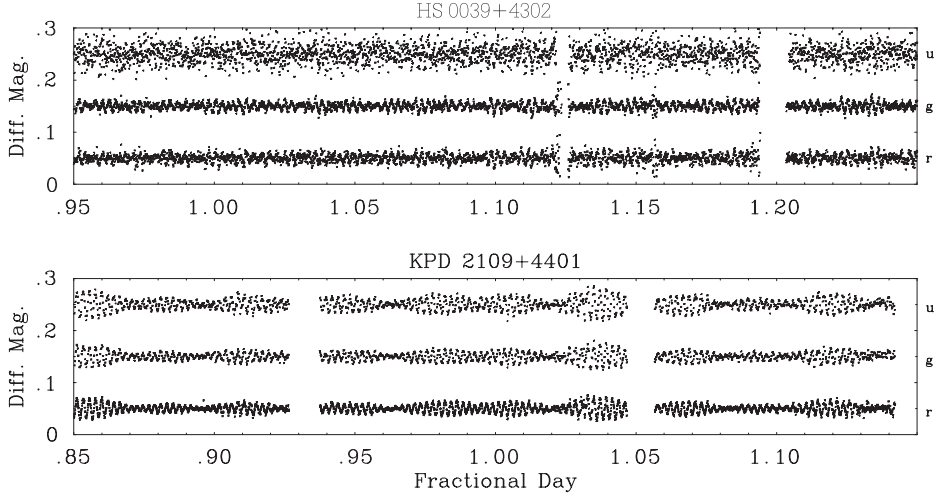


Figure 4. Partial ULTRACAM light curves in u' , g' and r' for sdBVs KPD 2109+4401 and HS 0039+3202 (adapted from Jeffery *et al.* 2004). Gaps are due to cloud.

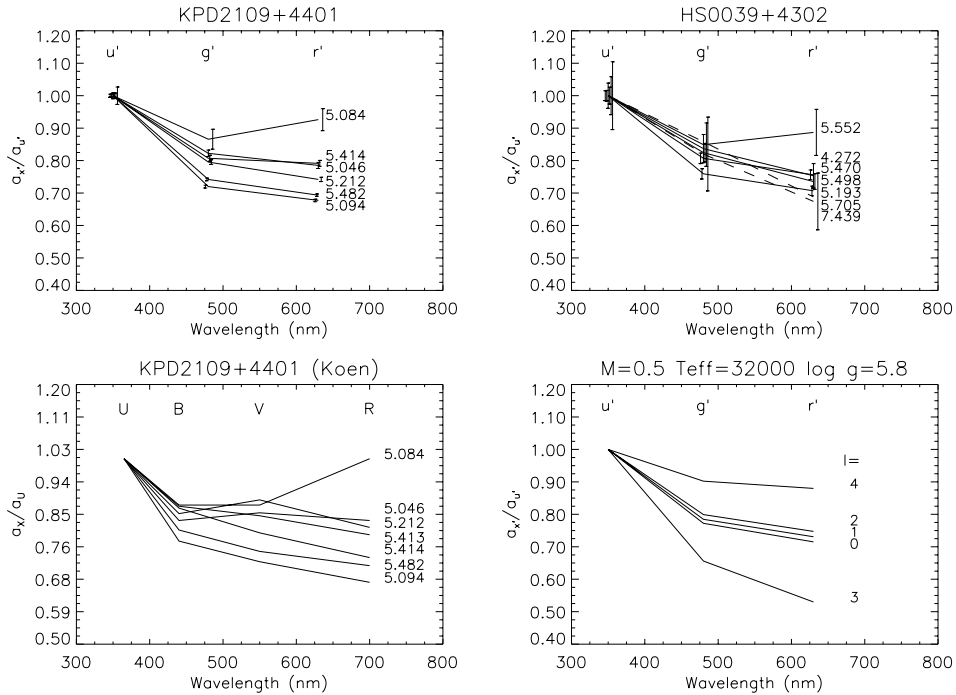


Figure 5. Amplitude ratios $a_{x'}/a_{u'}$ are shown for both KPD 2109+4401 and HS 0039+4302. The frequencies for each mode (mHz) are shown on the right. Modes in which $a_{u'} < 1.4$ mmag are shown as dashed lines. Error bars for each value are shown with small horizontal offsets for clarity. Earlier photometry of KPD 2109+4401 (a_x/a_U ; Koen 1998) is shown for comparison. The lower right panel shows theoretical colour amplitude ratios for a model with $M = 0.5 M_{\odot}$, $T_{\text{eff}} = 32000$ K, $\log g = 5.8$, $f = 5.4$ mHz and for $l = 0, \dots, 4$ from Ramachandran *et al.* (2004). Figure from Jeffery *et al.* (2004).

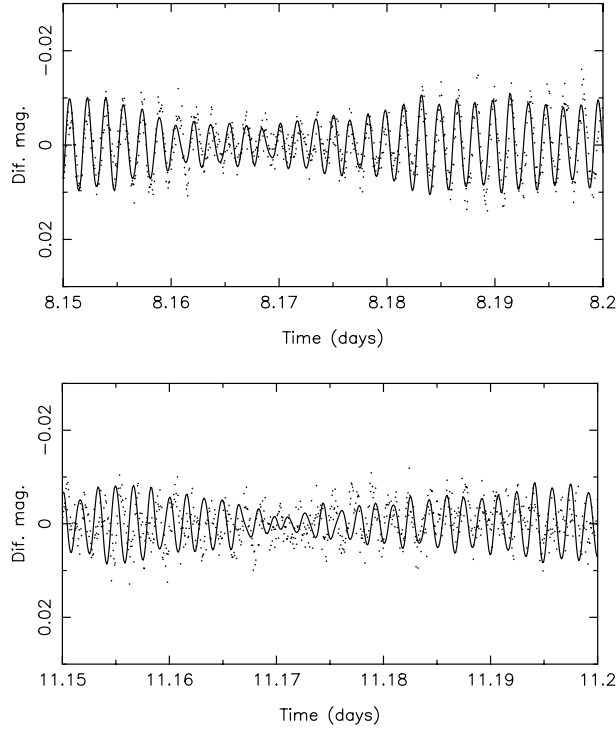


Figure 6. Two sections of the white light curve of PG 0014+067 together with the 19-frequency solution. Time (t) is JD-2453230. The change of amplitude of the dominant mode on a timescale of days is clearly visible. This is probably a consequence of rotational splitting of the dominant mode. This figure is reproduced from Jeffery *et al.* (2005).

in Fig. 4. From amplitude spectra obtained with these light curves, the amplitudes of several modes were measured in each of the three ULTRACAM channels, being u' , g' and r' of the SDSS system. The ratios $a_{g'}/a_{u'}$ and $a_{r'}/a_{u'}$ were computed and are shown in Fig. 5. Meanwhile, theoretical models of the colour amplitude ratios expected for non-radial oscillations of different spherical degree in sdBVs were computed by Ramachandran *et al.* (2004). From Fig. 5, it was clear that at least one oscillation in each target had to be a relatively high-degree ($l = 4$) mode. Predicted ratios for low-degree modes ($l = 0, 1, 2$) lie close together, so it is not possible to identify the degree of the observed modes so easily. Assuming the observed frequencies belong to modes with unique k, l values (*i.e.*, there is no rotational splitting), an additional constraint can be used. For a given degree l , modes of successive radial order k must be well-spaced in frequency, so modes of similar frequency cannot have the same l . By imposing such a constraint and by comparing the colour-amplitude ratios, it is possible to assign k and l values with some confidence. It is then possible to compare the observed frequency spectrum with theoretical models for non-radial oscillations in extended horizontal branch stars (cf. Charpinet *et al.* 2002), and hence to select which models best represent the star observed in terms of total mass, envelope mass, age, and other characteristics. A full description of the observations and the identification of the modes in these two stars is given by Jeffery *et al.* (2004).

5.2 2004: PG0014+067

A second ULTRACAM run to observe pulsating sdBVs was executed in September 2004. The target was the magnitude 16.5 star PG0014+067, previously the subject of a successful asteroseismological study by Brassard *et al.* (2001). This star shows ~ 20 independent frequencies in its light curve with frequencies between 5.7 and 12.9 mHz. Our aim was to check mode identifications deduced from the theoretical model that best fitted the frequency spectrum, and also to support a subsequent Whole Earth Telescope campaign to improve the overall frequency resolution. Although a much more challenging target than the 2002 targets, it was possible to observe this object with the WHT for ~ 6 hours on each of 6 successive nights. The results were less conclusive than before, but it was possible to show that the two dominant modes must be $l = 0, 1$ or 2, and to find evidence that the rotational period should be nearer to 4d than to the 1.35d reported previously (Fig. 6). Most major frequencies identified previously were confirmed, although cycle $d - 1$ aliases still contribute some uncertainties. Full details will be published by Jeffery *et al.* (2005).

6. Conclusions

We have shown how asteroseismology offers excellent prospects for exploring the internal structure and evolutionary status of subdwarf B stars. With ULTRACAM, we can obtain outstanding 3-channel light curves for pulsating sdBs down to 16th mag. These amplitudes are measurable to precisions < 0.5 millimags. and have enabled us to detect new frequencies. We have been able to identify l -values for modes with l up to 4 by ranking colour amplitude ratios and we have been able to assign the radial order n by demanding a realistic cadence from nodes of the same l . Future observations will exploit these techniques in more intensive studies of individual stars and hence provide data on the masses of sdBs and the chemical structure of their outer layers. We are just beginning to exploit the potential of longer-period sdBVs as probes of the deep interiors of extended horizontal branch stars.

There are substantial opportunities for ARIES astronomers to carry out high-speed photometry to make discovery observations and to participate in multisite campaigns, particularly to resolve periods for the long-period sdBVs.

Acknowledgements

The author is grateful to the NI DCAL and the UK PPARC for financial support, and to the ULTRACAM, WHT and AAT staff for practical support. He expresses particular thanks to ARIES for travel support to attend this conference.

References

- Brown, T. M., Bowers, C. W., Kimple, R. A. *et al.* 2000, *ApJ*, **532**, 308.
- Brown, T. M., Sweigart A. V., Lanz, T., Landsman W. B., Hubeny, I. 2001, *ApJ*, **562**, 368.
- Allard, F., Wesemael, F., Fontaine, G., Bergeron, P., Lamontagne, R. 1994, *AJ*, **107**, 1565.
- Aznar Cuadrado, R., Jeffery, C. S. 2001, *A&A*, **368**, 994.
- Billères, M., Fontaine, G., Brassard, P., Charpinet, S., Liebert, J., Saffer, R. A., Vauclair, G. 1997, *ApJ*, **487**, L81.
- Billères, M., Fontaine, G., Brassard, P., Charpinet, S., Liebert, J., Saffer, R. A., Bergeron, P., Vauclair, G. 1998, *ApJ*, **494**, L75.

- Brassard, P., Fontaine, G., Billères *et al.* 2001, *ApJ*, **563**, 1013.
- Charpinet, S., Fontaine, G., Brassard, P., Dorman, B. 1996, *ApJ*, **471**, L103.
- Charpinet, S., Fontaine, G., Brassard, P., Chayer, P., Rogers, F. J., Iglesias, C. A., Dorman, B. 1997, *ApJ*, **483**, L123.
- Charpinet, S., Fontaine, G., Brassard, P. 2001, *PASP*, 775.
- Charpinet, S., Fontaine, G., Brassard, P., Dorman, B. 2002, *ApJS*, **140**, 469.
- Green, E. M., Fontaine, G., Reed M. *et al.* 2003, *ApJ*, **583**, 31.
- Greenstein, J. L., Sargent, A. I. 1974, *ApJS*, **28**, 157.
- Gautschi, A., Saio, H. 1995, *Ann Rev. Astr. Astrophys.*, **33**, 75.
- Gautschi, A., Saio, H. 1996, *Ann Rev. Astr. Astrophys.*, **34**, 551.
- Heber, U. 1986, *AA*, **155**, 33.
- Heber, U., Hunger, K., Jonas, G., Kudritzki, R. P. 1984, *AA*, **130**, 119.
- Heber, U., Kudritzki, R. P., Caloi, V., Castellani, V., Danziger, J., Gilmozzi, R. 1986, *AA*, **162**, 171.
- Heber, U., Dreizler, S., Schuh, S. L. *et al.* 2003, In “White Dwarfs”, NATO Science Series II/105, p. 105.
- Heynderickx, D., Waelkens, C., Smeyers, P. 1994, *A&AS*, **105**, 447.
- Iben, I. Jr., Tutukov, A. V. 1985, *ApJS*, **58**, 661.
- Iben, I. Jr., Tutukov, A. V. 1986, *ApJ*, **311**, 753.
- Iben, I. Jr. 1990, *ApJ*, **353**, 215.
- Jeffery, C. S., Pollacco, D. 2000, *MNRAS*, **318**, 974.
- Jeffery, C. S., Dhillon, V. S., Marsh, T. R., Ramachandran, B. 2004, *MNRAS*, **352**, 699.
- Jeffery, C. S., Aerts, C., Dhillon, V. S., Marsh, T. R., Gänsicke, B. 2005, *MNRAS*, in press.
- Kilkenny, D. 2002, *ASP Conf. Series*, **259**, 356.
- Kilkenny, D., Koen, C., O’Donoghue, D., Stobie, R. S. 1997, *MNRAS*, **285**, 640.
- Kilkenny, D., Koen, C., O’Donoghue, D., van Wyk, F., Larson, K. A., Shobbrook, R., Sullivan, D. J., Burleigh, M. R., Dobbie, P. D., Kawaler, S. D. 1999, *MNRAS*, **303**, 525.
- Koen, C. 1998, *MNRAS*, **300**, 567.
- Koen, C., Kilkenny, D., O’Donoghue, D., van Wyk, F., Stobie, R. S. 1997, *MNRAS*, **285**, 645.
- Koen, C., O’Donoghue, D., Kilkenny, D., Lynas-Gray, A. E., Marang, F., van Wyk, F. 1998, *MNRAS*, **296**, 317.
- Maxted, P. F. L., Marsh, T. R., North, R. C. 2000, *MNRAS*, **317**, L41.
- Michaud, G., Bergeron, P., Wesemael, F., Heber, U. 1989, *ApJ*, **338**, 417.
- O’Donoghue, D. 1999, *ASP Conf. Series*, **169**, 149.
- O’Donoghue, D., Koen, C., Solheim, J.-E., Barstow, M. A., Dobbie, P. D., O’Brien, M. S., Clemens, J. C., Sullivan, D. J., Kawaler, S. D. 1998, *MNRAS*, **296**, 296.
- O’Toole, S. J. 2004, *A&A*, **423**, 250.
- O’Toole, S. J., Teixeira, T. C., Bedding, T. R., Kjeldsen, H. 2000, *ApJL*, **537**, L53.
- O’Toole, S. J., Falter, S., Heber, U. *et al.* 2004, *Ap&SS*, **291**, 457.
- Ramachandran, B., Jeffery, C. S., Townsend, R. H. D. 2004, *A&A*, **428**, 209.
- Reid, A., Bolton, C. T., Crowe, R. A., Fieldus, M. S., Fullerton, A. W., Gies, D. R., Howarth, I. D., McDavid, D., Prinja, R. K., Smith, K. C. 1993, *ApJ*, **417**, 320.
- Rutten, R., Gribbin, F., Ives, D., Bennett, T., Dhillon, V. 1997, “Drift-Mode CCD readout – User’s Manual”, Isaac Newton Group, La Palma.
- Saio, H. 1996, *ASP Conf. Series*, **96**, 361.
- Saio, H., Jeffery, C. S. 2000, *MNRAS*, **313**, 671.
- Stathakis, R. A., Johnston, H. M. 1997, AAO UM-2.3: The RGO Spectrograph Manual Version 3.1. Anglo-Australian Observatory, Epping.
- Telting, J. H., Aerts, C., Mathias, P. 1997, *A&A*, **322**, 493.
- Vauclair, G. 1996, *ASP Conf. Series*, **96**, 397.
- Wolf, V. M., Jeffery, C. S., Pollacco, D. 2002, *MNRAS*, **329**, 497.

Lessons for Asteroseismology from White Dwarf Stars

Travis S. Metcalfe

High Altitude Observatory, NCAR, P.O. Box 3000, Boulder CO 80307, USA.

e-mail: travis@hao.ucar.edu

Abstract. The interpretation of pulsation data for sun-like stars is currently facing challenges quite similar to those faced by white dwarf modelers ten years ago. The observational requirements for uninterrupted long-term monitoring are beginning to be satisfied by successful multi-site campaigns and dedicated satellite missions. But exploration of the most important physical parameters in theoretical models has been fairly limited, making it difficult to establish a detailed best-fit model for a particular set of oscillation frequencies. I review the past development and the current state of white dwarf asteroseismology, with an emphasis on what this can tell us about the road to success for asteroseismology of other types of stars.

Key words. Stars: interiors—stars: oscillations—white dwarfs.

1. Introduction

For the past six years, I have been trying to improve the way that we match theoretical models to observations of pulsating white dwarfs. I have made some progress transforming the art of model-fitting from a hands-on procedure to something more objective, more global, and more automated. The impact of this approach on the analysis of pulsating white dwarfs over the past few years suggests that it might also be able to improve our seismological modeling of other types of stars.

I am now interested in applying this same approach to the relatively recent (and much more difficult) observations of solar-like oscillations in other stars. Reading through the recent literature on this subject, I have been impressed by how the current state of model-fitting for these stars resembles the state of white dwarf modeling ten years ago. With this in mind, my goal here is to share some of the lessons I've learned while working on white dwarfs, with the hope that they will be useful for the future analysis of sun-like stars.

Before I go any further, I'd like to make it clear that I do not think the current analysis methods for sun-like stars are *wrong*. I do think that the current methods are limited in various ways, and these limitations may prevent us from learning all that we can learn from the observations. I hope to demonstrate this possibility with some anecdotes about our recent progress modeling the pulsating white dwarfs.

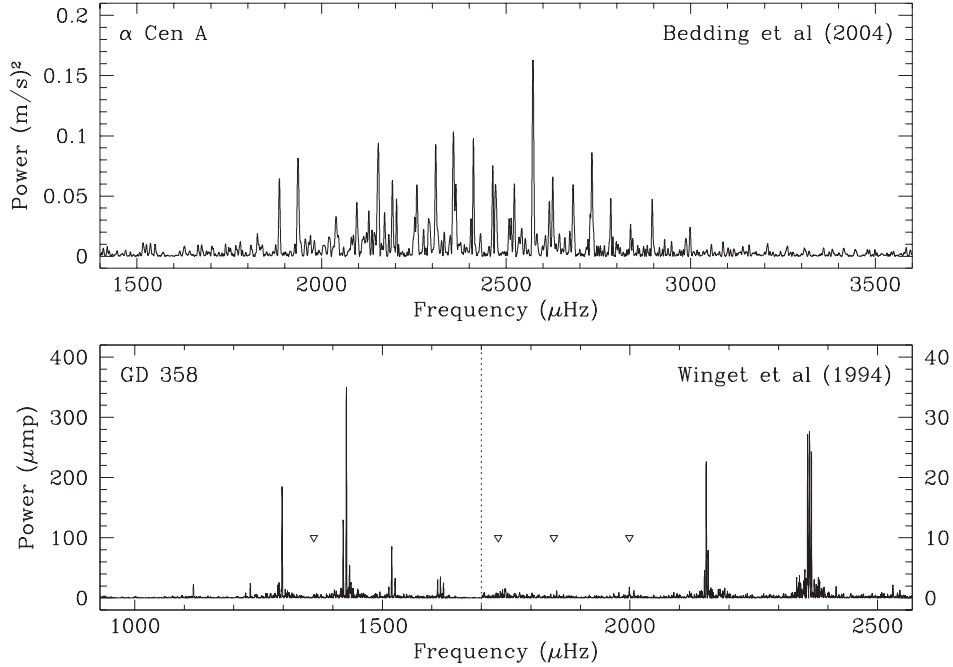


Figure 1. Power spectra of time-series measurements to detect solar-like oscillations in α Cen A published recently (top), compared to the pulsations observed in the white dwarf GD 358 published ten years earlier (bottom). The dotted line in the bottom panel divides the regions where the left and right vertical scales are applicable, and the open triangles mark the frequencies for several of the lower amplitude modes.

2. Similarities and differences

In case you haven't noticed, the Hertzsprung–Russell diagram is absolutely filled with pulsating stars (see Christensen-Dalsgaard 1998, Fig. 1). We believe that solar-like oscillations should be excited in just about any star with a convection zone. The basic idea is that convection generates sound waves at all frequencies (“white noise”, like static on a radio). Some of these frequencies are resonant inside the spherical cavity of the star and become long-lived global pulsations (standing waves). Such pulsations have recently been detected in several main-sequence, subgiant, and giant stars. The oscillation amplitudes and the frequency of maximum power in these stars agree reasonably well with our theoretical expectations (see Bedding & Kjeldsen 2003 for a recent review).

There are several classes of pulsating white dwarfs, almost equally spaced in $\log T_{\text{eff}}$. In each case, the pulsations seem to be excited by a similar mechanism (extra opacity from a near-surface partial ionization zone) but the surface chemical composition can be hydrogen, helium, or something else. Compared to other types of stars, white dwarfs are relatively simple from a physical perspective. Nuclear fusion is no longer active, and potentially complicating factors like fast rotation or strong magnetic fields are not a problem for most of these stars (Winget 1998).

The spherical symmetry of gravity allows us to describe stellar pulsations with spherical harmonic functions, the same mathematics that we use to describe electron

orbitals in quantum mechanics. We can assign each pulsation mode three quantum numbers: an l and m value to describe the pattern on the visible surface, and an n value (sometimes called k) for different radial overtones. For observations of stars, only modes with relatively low values of l can be detected, because the many bright and dark areas tend to cancel each other out for higher- l modes.

Asteroseismology of white dwarfs has advanced more quickly than that for other types of stars, in part because the pulsations are much less subtle. The total observed light variation is typically 10% or more, on very convenient timescales of 5–15 minutes. The light variations caused by solar-like oscillations are considerably smaller, so scintillation in the Earth's atmosphere severely limits our ability to detect them. However, the pulsations also cause variations in the spectral line profiles and equivalent widths, allowing us to detect them from the ground with time-series spectroscopic measurements.

From an observational standpoint, Fig. 1 illustrates just how similar the current status of sun-like stars is to the situation we had for white dwarfs ten years ago. In both panels we see a well-defined range of excited frequencies. For sun-like stars we see a series of almost equally spaced frequencies, which is characteristic of pressure (p -) modes. For white dwarfs the peaks are unevenly spaced in frequency, but evenly spaced in period. This is characteristic of the gravity (g -) modes.

From a theoretical standpoint, the observed pulsation periods are determined by the values of the sound speed (or Lamb frequency, L) and the buoyancy frequency (N) from the center of the star to the surface. The p -modes are excited at frequencies higher than both L and N , while the g -modes are excited at frequencies lower than both L and N (Unno *et al.* 1989). In a chemically uniform stellar model the pulsation modes are evenly spaced, but chemical stratification and variations in other relevant physical quantities cause discontinuities in L and N that lead to unevenly spaced modes. These deviations from uniform spacing (in frequency or period) contain detailed information about the interior structure of the star.

A fundamental difference between the observed modes in sun-like stars and white dwarfs is that we generally see modes of only one spherical degree in white dwarfs (mostly $l = 1$, but in some cases $l = 2$). In sun-like stars, we usually see radial ($l = 0$) and non-radial ($l = 1, 2$ and sometimes 3) modes excited simultaneously. Consequently, there is a historical difference in the way the pulsation data are usually presented. In Fig. 2, this difference has been circumvented by rotating the white dwarf plot. This allows us to see more readily how the current state of model-fitting for sun-like stars is also very similar to what was being done with white dwarfs ten years ago.

In white dwarf models, the mean period spacing is related to the total stellar mass, while the deviations from the mean period spacing (dP) tell us about the thickness of the stratified surface layers (Bradley *et al.* 1993). In models of sun-like stars, the large frequency spacing ($\Delta\nu$) is related to the mean density of the star. The deviations that split modes with odd or even spherical degree (the so-called small frequency separation, $\delta\nu$) tell us about chemical gradients in the interior (see Brown & Gilliland 1994). The current level of sophistication for model-fitting of sun-like stars allows us to match the overall patterns of the oscillations, but there is room for improvement in the detailed fit to the individual modes. This is just where white dwarf modeling was in 1994.

The story of how we improved our approach to white dwarf modeling contains some important lessons for the modeling of sun-like stars (see also Metcalfe 2003a).

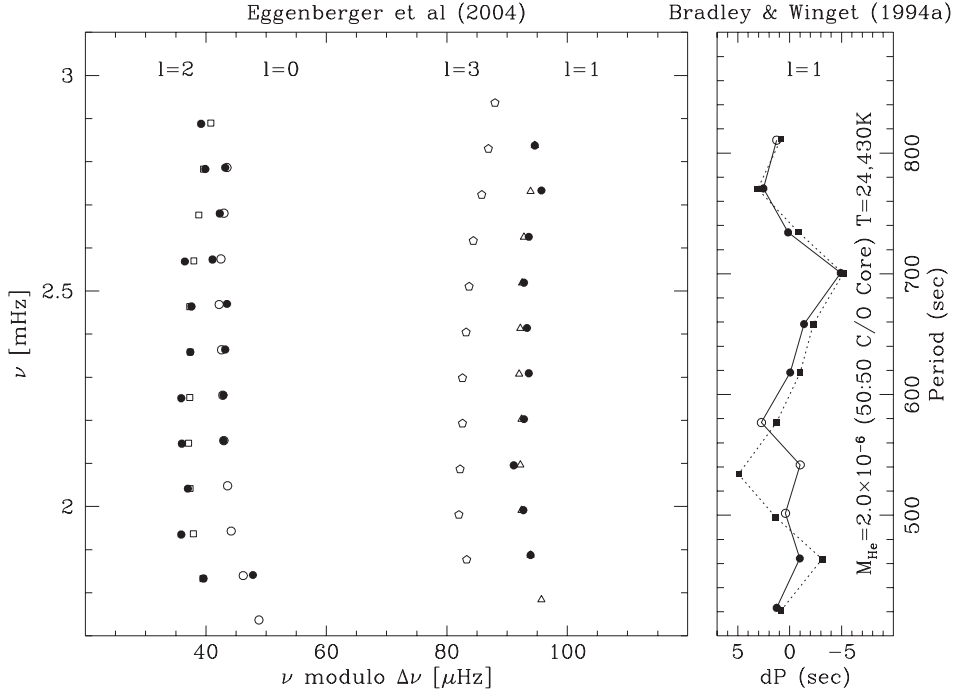


Figure 2. A recent diagram of the frequencies observed in α Cen A plotted against their deviations from the large frequency spacing (left), and a similar diagram from ten years earlier showing the periods observed in GD 358 plotted against their deviations from the mean period spacing (right). In both cases the theoretical models match the overall patterns, but some improvement is still needed to match the fine details.

This is particularly relevant now, since the next few years promise to unleash a flood of new observations of solar-like oscillations in other stars. We have already seen what is possible with a relatively modest space mission devoted to asteroseismology, the Canadian MOST satellite (Walker *et al.* 2003). Forthcoming missions with larger apertures, such as COROT (Baglin *et al.* 2002) and Kepler (Borucki *et al.* 2003) are likely to extend these successes and bring new surprises. Unfortunately, none of these missions will contribute new data on the faintest pulsating stars, so future advances for white dwarfs and variable sdB stars will most likely continue to come from ground-based observations.

3. Using observational constraints

In the current model-fitting approach for sun-like stars, it is common to use external observational constraints on the luminosity and effective temperature to restrict the search to a narrow range of model parameters (e.g., see Di Mauro *et al.* 2003, Fig. 1). This guarantees that the seismological model will be consistent with these constraints, but it eliminates the advantage of the pulsation data. Even if we find a reasonable fit to the pulsation modes, how do we know it couldn't be improved if we relaxed or removed the constraints? The power of asteroseismology is that it provides *independent* constraints on the stellar structure. If the stellar models are accurate representations

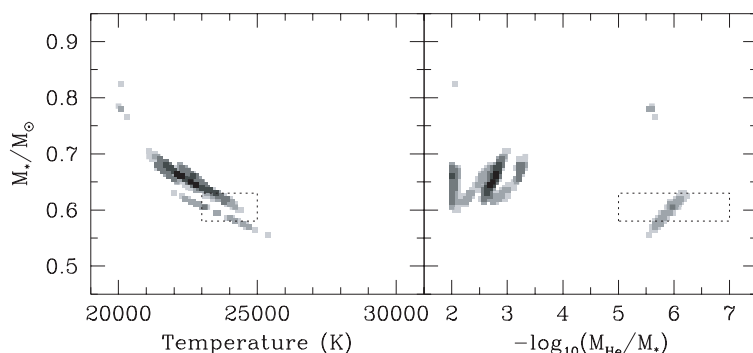


Figure 3. By searching a much broader range of parameter values compared to previous studies (dotted), we discovered a better match (darker) to the observations of GD 358. Our global search also revealed a strong sensitivity of the pulsations to the core composition, which eventually led to precise constraints on an important nuclear reaction rate.

of the real stars, then the pulsation modes should lead to the “observed” luminosity and effective temperature by themselves. If they do, this lends additional credibility to the resulting model. If instead the seismological data lead to a best-fit model that *disagrees* with the external constraints, then something is missing from the model; and that’s exactly what we hope to learn through asteroseismology. We will never find a disagreement if we don’t look for it.

To illustrate the importance of defining the range of the search as broadly as possible, let me tell you about the helium-atmosphere (DBV) white dwarf GD 358. The observations of this star came from the Whole Earth Telescope (Nather *et al.* 1990; Winget *et al.* 1994), and the initial asteroseismological analysis was done by Bradley & Winget (1994a). For our new approach using a parallel genetic algorithm (Metcalf & Charbonneau 2003), we decided to define the range of the search based only on the physics of the model whenever possible.

For DBV stars, the three most important parameters are the stellar mass, the effective temperature, and the thickness of the surface helium layer. For our initial study we searched masses between 0.45 and $0.95 M_{\odot}$. The observed mass distribution of white dwarfs is strongly peaked near $0.6 M_{\odot}$ (Napiwotzki *et al.* 1999). Lower mass white dwarfs are theoretically expected to have a helium core, and take longer than the age of the universe to form through single-star evolution. Higher mass white dwarfs are extremely rare. We searched effective temperatures between $20,000$ and $30,000$ K, which easily encompasses the full range of the DBV instability strip whether or not trace amounts of hydrogen are allowed in the envelopes (Beauchamp *et al.* 1999). We allowed the surface helium layer to be between 10^{-2} and $\sim 10^{-8}$ in fractional mass. Thicker layers would initiate nuclear burning at the base of the layer (ruled out by the observations), and for thinner layers our models do not pulsate (Bradley & Winget 1994b).

The full range of our search is shown in Fig. 3, along with the range of the original search in 1994 (dotted). We found the same best-fit model as Bradley & Winget when we restricted ourselves to their search range, but our broader search revealed an even better solution (darker) outside of the range they considered (Metcalf *et al.* 2000). This global approach also taught us that our models are much more sensitive to the core

composition than we had previously thought. This is significant because the oxygen mass fraction in the core of a white dwarf is largely determined by the relative rates of the 3α and the $^{12}\text{C}(\alpha, \gamma)^{16}\text{O}$ nuclear reactions during the red giant phase. The latter reaction is poorly constrained by laboratory data, which are extrapolated over six orders of magnitude in the cross-section to reach the energies relevant to red giant stars.

Motivated by our newly-discovered sensitivity to the core composition, we added adjustable C/O profiles to our models and performed a new global search. While our match to the observed pulsation modes improved slightly by considering a broader range of model parameters, optimizing the C/O profile improved the match dramatically (see Metcalfe *et al.* 2001, Fig. 4d). Since then we have applied the same method to another DBV white dwarf, CBS 114, leading to two independent constraints on the reaction rate that agree with each other (Metcalfe 2003b). The two values also agree with recent extrapolations from laboratory measurements (Angulo *et al.* 1999), but are considerably more precise.

By letting the pulsation modes speak for themselves we improved our fit to the observations, and in the process we discovered a previously unknown sensitivity in the models that ultimately led to constraints on an important nuclear reaction rate. Along the way, we found inconsistencies between some of the derived model parameters and other observations (e.g., spectroscopic masses and temperatures) that also prompted us to improve the model physics. None of this would have been possible if we had defined our search too narrowly.

4. Individual modes and spacing

Early attempts to detect solar-like oscillations in other stars produced Fourier spectra that did not clearly resolve individual frequencies, but showed excess power roughly where the excited modes were expected. Using auto-correlation techniques, these observations could infer the large frequency spacing, which could then be used to estimate the mean density of the star (see Kjeldsen *et al.* 1995, Fig. 3). More recent observations can now confidently resolve the individual modes, but auto-correlation is often still used to determine the large frequency spacing. Because the models may suffer from systematic errors, some recent model-fitting attempts have concentrated on the frequency spacing rather than matching the individual frequencies. This may be useful for a preliminary analysis, but the individual modes contain additional information that is lost by matching only the spacing. Using the individual modes can also produce clues about the source of any systematic errors, allowing us to improve the input physics of our models.

To illustrate the benefit of matching the individual modes, and not just the mode spacing, let me tell you about our experience with the hydrogen-atmosphere (DAV) white dwarf BPM 37093. When a white dwarf gets cool enough, the C/O core eventually undergoes a phase transition from liquid to solid. It crystallizes, from the center outward. This is important because the crystallization process releases latent heat, which delays the gradual cooling of the star. Since we can use the coolest white dwarfs in any stellar population (the Galactic disk and halo, open and globular clusters) to determine the age, it is crucial that we calibrate this cooling delay. Otherwise, we will underestimate the true ages.

A typical white dwarf with a mass of $0.6 M_{\odot}$ will begin to crystallize when it cools to an effective temperature between 6000 and 8000 K, depending on the core

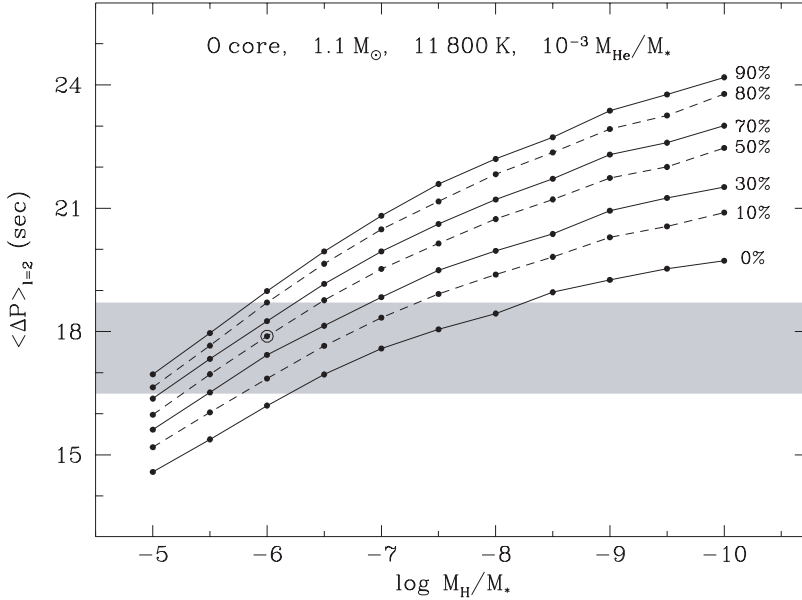


Figure 4. When our fit used only the *mode spacing* (shaded), we were unable to distinguish between thinner H layers and a larger crystallized mass fraction in BPM 37093 — both parameters increased the mode spacing. By exploiting the information in the *individual modes*, one model (circled) stands out as the best in this small grid. Of course, a simultaneous optimization of all of the parameters is needed to find the global solution.

composition. More massive white dwarfs have higher central pressures, so they will begin to crystallize at higher temperatures. BPM 37093 is massive enough ($\sim 1.1 M_\odot$) that it has theoretically started to crystallize while at an effective temperature inside the DAV instability strip ($\sim 12,000$ K). The pulsations do not penetrate the solid core, so the inner boundary for each mode is located at the edge of the crystallized core rather than near the center of the star. This changes the period of each mode, compared to what it would have been in the absence of crystallization. As a consequence, we can calibrate the theory of crystallization empirically by modeling the pulsations in this star.

The observations of BPM 37093 were obtained by the Whole Earth Telescope (Kanaan *et al.* 2000, 2005), and the initial theoretical study came from Montgomery & Winget (1999). This early analysis concentrated on the average period spacing of the modes, since the main effect of the crystallized core is to spread the pulsation periods farther apart. However, they found a troubling degeneracy between changes to the crystallized mass fraction and changes to the thickness of the surface hydrogen layer, leading to ambiguous results from the average period spacing (see Fig. 4). By exploiting the information contained in the *individual periods*, we were able to lift this degeneracy and determine a unique best-fit model. Of course, in the small grid of models shown in Fig. 4, we have fixed several other important parameters, so the true best-fit model requires a more global exploration.

This is a huge computational problem, since we need to consider the stellar mass, the effective temperature, the thickness of the helium mantle and the surface hydrogen layer, the core composition, and the crystallized mass fraction, all simultaneously. Our initial fits to the individual periods confirmed that BPM 37093 is substantially

crystallized (Metcalfe *et al.* 2004), but the search was too limited to determine the exact fraction. Our latest results look promising (Metcalfe *et al.* 2005), and have recently been confirmed by Brassard & Fontaine (2005). But this expanded search still required that we fix the mass and core composition, and there remain many locally optimal models, so our search for a global solution continues.

By using the individual modes, and not just the spacing, we lifted the degeneracy between the crystallized mass fraction and the surface hydrogen layer mass in models of the DAV white dwarf BPM 37093. This led to the first direct confirmation of crystallization theory, and opened the door to a more global exploration of the models which is still in progress. If we had confined ourselves to fitting the average period spacing, we would never have obtained this first glimpse inside of a crystallized star.

5. Summary

Let me summarize the main points.

- The future of asteroseismology for sun-like stars is primarily in space.
We soon expect many new observations from current and planned satellite missions. It will be a distinct advantage to have in place the analysis methods that will help us make sense of these data efficiently.
- *Lesson 1:* Let the pulsations speak for themselves.
To interpret these data, it is important that we avoid introducing our subjective biases by defining the range of our model-fitting search too narrowly. The asteroseismic models should agree with independent measurements without coercion.
- *Lesson 2:* Use the individual modes, not just the spacing.
The individual frequencies contain additional information that is lost if we focus only on the overall pattern of frequency spacing. This could unnecessarily limit what we are ultimately able to derive from the observations.
- The future of asteroseismology for fainter stars is multi-site.
Upcoming space missions will not affect the future of the fainter pulsators. Advances in our understanding of white dwarfs and variable subdwarf B stars will continue to come from multi-site campaigns, possibly with larger telescopes.

This is a very exciting time for asteroseismology. It is the beginning of a new era, when we will improve our knowledge of solar physics by studying sun-like stars. This will allow us to explore the past and future of our own sun by observing pulsations in solar analogues, calibrating our models at different evolutionary stages and under various physical conditions. For the white dwarf stars, we are closer to the end of an era. Over the past decade we have steadily improved our models and our analysis methods, charting a course that is just beginning for sun-like stars. If we are clever enough, we can use what we have learned along the way to accelerate our progress on other types of stars.

Acknowledgements

This work was supported by the National Science Foundation through an Astronomy & Astrophysics Post-doctoral Fellowship (AST-0401441). I would like to thank Jørgen

Christensen-Dalsgaard for motivating me to turn my attention to sun-like stars, and Tim Brown for helping to bring this project to the High Altitude Observatory at the National Center for Atmospheric Research.

References

- Angulo, C. *et al.* 1999, *Nucl. Phys. A*, **656**, 3.
 Baglin, A. *et al.* 2002, *Astro. Soc. Pacif. Conf.*, **259**, 626.
 Beauchamp, A. *et al.* 1999, *Astrophys. J.*, **516**, 887.
 Bedding, T. R., Kjeldsen, H. 2003, *Pub. Astro. Soc. Aust.*, **20**, 203.
 Bedding, T. R. *et al.* 2004, *Astrophys. J.*, **614**, 380.
 Borucki, W. J. *et al.* 2003, *Astro. Soc. Pacif. Conf.*, **294**, 427.
 Bradley, P. A., Winget, D. E., Wood, M. A. 1993, *Astrophys. J.*, **406**, 661.
 Bradley, P. A., Winget, D. E. 1994a, *Astrophys. J.*, **430**, 850.
 Bradley, P. A., Winget, D. E. 1994b, *Astrophys. J.*, **421**, 236.
 Brassard, P., Fontaine, G. 2005, *Astrophys. J.*, **622**, 572.
 Brown, T. M., Gilliland, R. L. 1994, *Ann. Rev. Astron. Astrophys.*, **32**, 37.
 Christensen-Dalsgaard, J. 1998, *Astrophys. Space Sci.*, **261**, 1.
 Di Mauro, M. P. *et al.* 2003, *Astron. Astrophys.*, **404**, 341.
 Eggenberger, P. *et al.* 2004, *Astron. Astrophys.*, **417**, 235.
 Kanaan, A. *et al.* 2000, *Baltic Astron.*, **9**, 87.
 Kanaan, A. *et al.* 2005, *Astron. Astrophys.*, **432**, 219.
 Kjeldsen, H. *et al.* 1995, *Astron. J.*, **109**, 1313.
 Metcalfe, T. S., Nather, R. E., Winget, D. E. 2000, *Astrophys. J.*, **545**, 974.
 Metcalfe, T. S., Winget, D. E., Charbonneau, P. 2001, *Astrophys. J.*, **557**, 1021.
 Metcalfe, T. S., Charbonneau, P. 2003, *J. Comput. Phys.*, **185**, 176.
 Metcalfe, T. S. 2003a, *Astrophys. Space Sci.*, **284**, 141.
 Metcalfe, T. S. 2003b, *Astrophys. J.*, **587**, L43.
 Metcalfe, T. S., Montgomery, M. H., Kanaan, A. 2004, *Astrophys. J.*, **605**, L133.
 Metcalfe, T. S., Montgomery, M. H., Kanaan, A. 2005, *Astro. Soc. Pacif. Conf.*, **334**, 465.
 Montgomery, M. H., Winget, D. E. 1999, *Astrophys. J.*, **526**, 976.
 Napiwotzki, R., Green, P. J., Saffer, R. A. 1999, *Astrophys. J.*, **517**, 399.
 Nather, R. E. *et al.* 1990, *Astrophys. J.*, **361**, 309.
 Unno, W. *et al.* 1989, *Nonradial oscillations of stars*, 2nd edn, Univ. Tokyo Press.
 Walker, G. *et al.* 2003, *Pub. Astro. Soc. Pacif.*, **115**, 1023.
 Winget, D. E. *et al.* 1994, *Astrophys. J.*, **430**, 839.
 Winget, D. E. 1998, *J. Phys. Cond. Matt.*, **10**, 11247.

On the Analysis of Light Curves in Asteroseismology

David L. Mary

Aryabhata Research Institute of Observational Sciences, Manora Peak, Nainital 263 129, India.

e-mail: dmary@upso.ernet.in

Abstract. We provide a detailed introduction to the main problems arising when analyzing light curves in asteroseismology. Attention is first paid to the signal model delivered by the pulsating stars and to the noise sources corrupting this model in photometric observations. The main pitfalls and ambiguities occurring in Fourier analysis are summarized and illustrated. Some classical, Least Squares (LS) based methods for spectrum analysis are analyzed and commented on from the point of view of ill-posed problems. The insight that can be gained from such analyses is discussed.

Key words. Fourier analysis—noise sources—ill-posed problems.

1. Introduction

Pulsating stars form a subclass of variable stars. The luminosity variations exhibited by these stars depend upon their mass, structure and chemical composition. The purpose of asteroseismology is to infer physical parameters for the stars (internal structure, effective temperature, magnetic field, gravity, mass, inclination angle, . . .) by analyzing their pulsation patterns. The overall strategy can be divided in to two main parts. In the first part, the data are collected under the form of variability light curves. After a series of operations on these data (reduction, noise analysis, spectrum estimation, see S. Joshi's paper in this issue for examples of roAp stars), the eigenfrequencies are determined, leading to the eigenmodes of pulsation. The second part deals with designing numerical models of stars (and of the corresponding pulsation patterns) by perturbing fluid equations (see e.g., Finley *et al.* (1997) for atmosphere's models). The better the agreement with the data, the better the model (see e.g., Castanheira *et al.* (2004) for a detailed example of a White Dwarf).

The present paper introduces in detail the first part of the above strategy: we present an overview of the techniques and problems related to the determination of the pulsation's frequencies. In section 2, attention is first paid to the mathematical model for signals obtained from pulsating stars; important pitfalls and ambiguities occurring in Fourier analysis are illustrated. We then turn to the noise sources corrupting this model in photometric observations. The third section comments and analyzes two classical methods (CLEAN, SPD) for Fourier Transform (FT) analysis. General Least Squares (LS) based methods are further analyzed and commented from the point of view of ill-posed problems in section 4. We discuss how this point of view can improve the analysis of light curves and lead to different, sometimes more appropriated analysis methods. Section 5 draws some conclusions.

2. Light curves in asteroseismology

2.1 Signal model, Fourier analysis and data gaps

Many pulsating stars exhibit small amplitude variability (e.g., sun, rapidly oscillating A peculiar (roAp) stars, white dwarfs, ...). Small amplitude oscillations of spherical objects can be described in terms of spherical harmonics. For a particular mode, the oscillations of scalar quantities such as radial velocity or luminosity can be written in the form: $m(t) = m_0 + a_0 \cos(2\pi \nu_0 t - \phi_0)$, where $m(t)$ and m_0 are respectively the instantaneous and mean scalar quantity (magnitude in the following), a_0 , ν_0 and ϕ_0 are respectively the amplitude, frequency and phase of the mode. Several modes may be simultaneously excited in the star, in which case the variable magnitude becomes $m(t) = m_0 + \sum_i a_i \cos(2\pi \nu_i t - \phi_i)$.

For larger (but still periodic) amplitude oscillations, the elementary variability shape may not be sinusoidal anymore (e.g., Cepheids). In this case, peaks appear in the FT at multiples of the fundamental frequency (1 over main period). Hence the signal model is again of the form $m(t) = m_0 + \sum_i a_i \cos(2\pi \nu_i t - \phi_i)$, where a_i , ν_i and ϕ_i are respectively the amplitude, frequency and phase of the harmonics. The respective $\{a_i\}$ and $\{\phi_i\}$ determine the shape of the curve (see e.g., Poretti 2002). In all cases, the variability signal above is observed during a finite time. Assuming that no noise corrupts the light from the star to the detector, the observed magnitude becomes: $s(t) = m(t) \times w(t)$, where $w(t)$ is usually referred to as *observing* or *temporal window*.

The main justification of FT analysis is (in principle) to allow one to isolate the elementary frequencies of the signal. This is useful since several modes/harmonics are in general present in variability curves, resulting in (sometimes very) complicated spectra. Dropping the constant term m_0 in the signal $s(t)$ above, the FT can be written as

$$\begin{aligned} \widehat{s}(\nu) &= \widehat{m}(\nu) \star \widehat{w}(\nu) \\ &= \frac{1}{2} \sum_{i=1}^K a_i e^{j(\phi_i - \frac{\pi}{2})} \widehat{w}(\nu - \nu_i) - a_i e^{-j(\phi_i - \frac{\pi}{2})} \widehat{w}(\nu + \nu_i), \end{aligned} \quad (1)$$

where $\widehat{}$ and $\star$ denote respectively FT and convolution, $j = \sqrt{-1}$ and $\widehat{w}(\nu)$ is the *spectral window*. Equation (1) shows that we face a deconvolution problem: from $\widehat{s}(\nu)$, we seek to reconstruct $\widehat{m}(\nu)$.

Below are some examples.

- (a) The simplest case is that of one single mode (with parameters a_0, ν_0, ϕ_0) observed during one single observation night. Instead of a Dirac delta function, the contribution of the spectral window appears at the frequency $\nu_0 = 1/T_0$. For box-car observing window, $\widehat{w}$ is a sinc function ($\text{sinc } x = \sin x/x$). The power spectrum (PS, or periodogram, squared FT modulus) becomes: $|\widehat{s}(\nu)|^2 = \frac{1}{4} T^2 a_0^2 \text{sinc}^2[\pi T(\nu - \nu_0)]$, where we assumed that the observation duration T is much greater than the oscillation period T_0 . The *frequency resolution* (minimum frequency spacing for two different frequencies to be accurately localized) corresponds basically to the width of the sinc function ($2/T \approx 70 \mu\text{Hz}$ for one 8-hour night). Hence, the longer the observation, the better the resolution.
- (b) In the case of two oscillations (with parameters $\{a_1, \nu_1, \phi_1\}$ and $\{a_2, \nu_2, \phi_2\}$) observed during one clear night, the PS becomes

$$\begin{aligned}
|\hat{s}(\nu)|^2 = & \frac{1}{4} T^2 \{a_1^2 \text{sinc}^2[\pi T(\nu - \nu_1)] + a_2^2 \text{sinc}^2[\pi T(\nu - \nu_2)]\} \\
& + 2a_1 a_2 \text{sinc}[\pi T(\nu - \nu_1)] \text{sinc}[\pi T(\nu - \nu_2)] \\
& \times \cos[\pi T(\nu_2 - \nu_1) - (\phi_2 - \phi_1)].
\end{aligned} \tag{2}$$

The PS is not just the superimposition of the two elementary power spectral responses of example (a) because of the third (interference) term. For a given T , this term is important whenever ν_1 and ν_2 are close and T is not sufficiently large w.r.t. the pulsation period. In this case, the phase difference is important as well. This is illustrated in Fig. 1, left panels, with $\nu_1 = 1$ mHz and $\Delta\nu = \nu_1 - \nu_2 = 42 \mu\text{Hz}$. The beating of the two frequencies creates an equivocal representation of the signal. The ambiguity arises in the variability light curve also, which appears as one single cosine amplitude modulated mode even though two closely separated modes are present. The situation of closely separated modes is frequent in asteroseismology (magnetic (Unno *et al.* 1989) and rotational (Ledoux 1951) splitting); the presence of amplitude modulated modes also (see Handler 2004 for a review of this phenomenon in different classes of pulsating stars). In the latter case, the more damped the oscillation, the less narrow the corresponding spectral representation (see Christensen-Dalsgaard (2003), p. 20; Samadi *et al.* (2003) and references therein for more on stellar oscillations' excitation and damping). In both cases unfortunately, the observer has no mean to determinate what is really happening in the star, apart from increasing the resolution. From the PS curves (Fig. 1, left panels), one sees that the maxima are *not* at the actual frequencies. Note also that frequencies only do not tell the whole story: the PS strongly depends on the phases of the oscillations. According to $\Delta\phi = \phi_2 - \phi_1$ in equation (2), the maxima may be moved either inside or outside the actual frequency pair.

- (c) A solution for increasing the resolution is to observe the same star during two (or more) consecutive nights. In this case, the PS becomes for a single oscillation $|\hat{s}(\nu)|^2 = T^2 a_0^2 \text{sinc}^2[\pi T(\nu - \nu_0)] \cos^2[\pi d(\nu - \nu_0)]$, where d is the time interval between the two observation sessions (typically $d = 24$ h). The sinc^2 spectral window's envelope gets modulated by a $\cos^2$ function, yielding a fine structure. The side lobes ("one-day" aliases) are separated from the main lobe by $11.5 \mu\text{Hz}$ typically ($\pm 1 d^{-1}$, $\pm 2 d^{-1}$, etc.).
- (d) Let us now take a look at the previous case of two oscillations (example (b)), observed during two nights (Fig. 1, right panels). On the one hand, the maxima are closer to the original frequencies than in the one-night case since the interference term's contribution has been reduced (T has been increased). On the other hand, nine main peaks appear now in the PS. The fine structure (one-day aliases) are clearly visible.

The effects described above complicate the frequency representation of the oscillations. They can be far more complicated to understand when $m(t)$ is unknown, and when noise corrupts the data (see Christensen-Dalsgaard & Gough 1982).

2.2 Noise sources corrupting the model: The case of photometric observations

The pulsation signal $m(t)$ of section 2.1 suffers first from reddening and absorption when propagating through the interstellar medium. The atmosphere presents sky

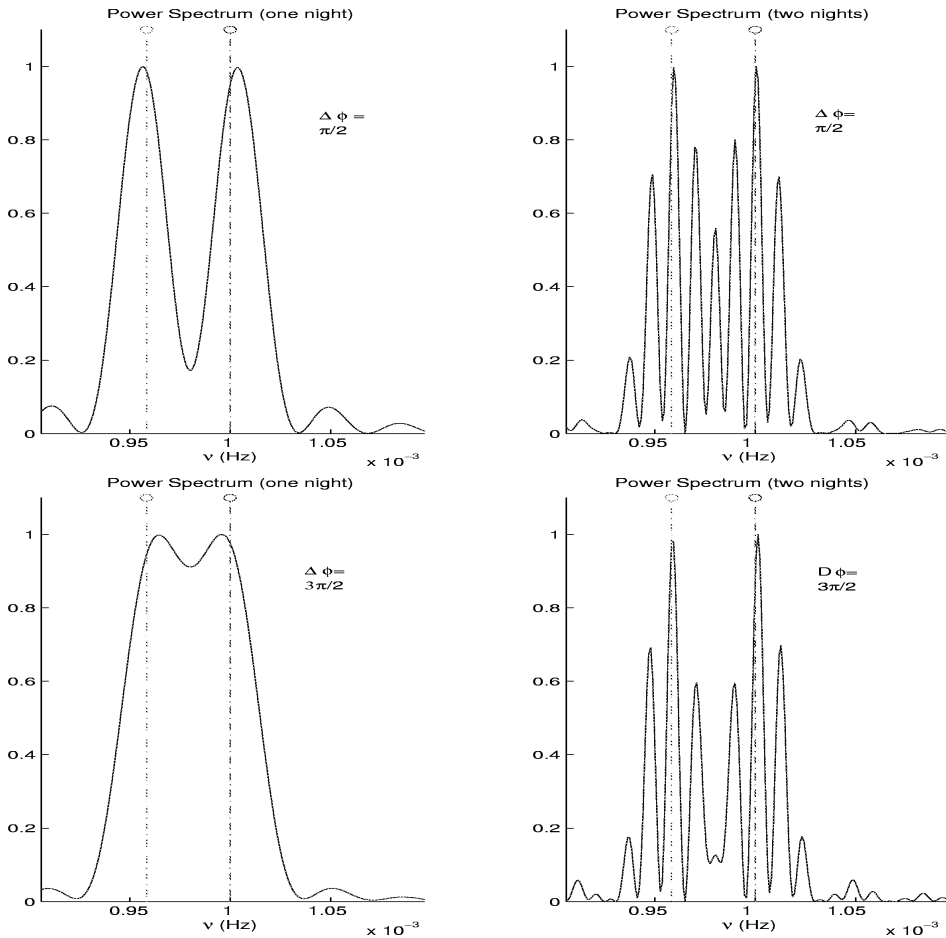


Figure 1. Data with gaps, two oscillations: comparison of the power spectra for one night (left panels) and two nights (right panels). Upper panels: $\Delta\phi = \pi/2$. Lower panels: $\Delta\phi = 3\pi/2$.

transparency variations depending on airmass, humidity, dust, etc.; this yields noise at frequencies less than ≈ 1 or 2 mHz. Atmosphere also produces *scintillation* caused by atmosphere's density fluctuations (Warner 1988), the corresponding noise's energy being at higher frequencies (Dravins *et al.* 1998). The former is usually removed by subtracting a few sinusoids to the data spectrum (*prewhitening method*, Ponman 1981). The latter is often larger than photon noise for relatively (e.g., roAp) bright stars. It can be decreased by using larger apertures (Young 1967) – and by observing from Antarctica, see Fossat's paper in this issue. Indeed, cosmic rays, clouds, planes, meteors, etc. may cause bad data points creating gaps. The signal that eventually falls on the detector is amplified according to the random gain and to the dead time of the photomultipliers (equivalent though less pronounced noises occur with CCDs) and the resulting signal is shaped for digitization. All these operations yield further distortions. Many other insidious noise sources may (and often do) occur: e.g., moisture/dust on optics and filters, telescope tracking oscillations (spurious periods of ≈ 2 to 4 sidereal minutes), moonlight reflection inside the telescope (spurious periods according to the

dome rotation), etc. (see Martinez 1993 for an extensive list). Consequently, it must be kept in mind that the signal model of section 2.1 is the very idealization of what the recorded “oscillation signal” may actually look like. The validity of the model may sometimes be highly questionable. On the other hand, the noise list above shows how difficult it can be to describe analytically departures from this model. Indeed, standard reduction techniques help reducing these effects, but to some extent only, and they may lead to artifacts as well (Balona 2002).

The comments above highlight how the simultaneous effects of gaps in the data, beating between close frequencies, damping and various noises can be confusing and lead to erroneous conclusions in the frequency analysis. If one goes for FT analysis, phase information is essential and amplitude rather than power spectra must be preferentially investigated. In order to increase the frequency resolution/reduce the data gaps, astronomers are often using multi-site, worldwide campaigns (Whole Earth Telescope, WET: e.g., Nather *et al.* 1990; STEllar PHotometry International: e.g., Michel *et al.* 1992; Delta Scuti Network: e.g., Breger *et al.* 1995). Because uncooperative weather, atmospheric and instrumental noises are unavoidable, the pitfalls and noises summarized in sections 2.1 and 2.2 remain major limiting factors for the accuracy of the variability analysis. As such, their nature and effects deserve to be as precisely understood as possible.

3. FT-based methods

Several attempts have been made to analyze noisy Fourier spectra in some automated and clever way. Two of them are discussed here: CLEAN is widely used — but often as a “black-box” unfortunately, as confessed by experienced asteroseismologists. The Several Peaks Deconvolution (SPD) method may be considered as a “one-shot” variant of CLEAN.

3.1 The CLEAN method

The CLEAN method was conceived by Jan Hogbom (1974) for eliminating side lobes in radio interferometry. Schwarz (1978) showed that CLEAN is equivalent to LS fitting. CLEAN assumes as signal model $s(t) = \sum_{i=1}^K a_i \sin(2\pi \nu_i t + \phi_i) \times w(t)$, so that the associated FT is as in equation (1). The algorithm performs the simple iterations (in the sequel, the symbol $\sim$ denotes estimates):

- (1) Find the frequency $\tilde{\nu}_i$, amplitude $\tilde{a}_i$ and phase $\tilde{\phi}_i$ of the largest peak,
- (2) Subtract the scaled contribution of $\hat{w}(\nu - \tilde{\nu}_i)$ from the FT (1),
- (3) Back to 1, or stop iterating as the FT residual is inferior to the noise level.

An example of CLEAN for a simulated oscillation signal is presented in Fig. 2 ($K = 5$). In this case, amplitudes, frequencies and phases are found with a relative precision of 10^{-3} . Note however that, even without noise, the residual is not zero. Indeed, at each iteration the maximum peak’s frequency $\tilde{\nu}_i = \nu_i + \delta\nu_i$ is displaced from the true frequency ν_i because of interferences. The estimated amplitude and phases are slightly incorrect as well (say $a_i + \delta a_i$ and $\phi_i + \delta\phi_i$). Hence, terms of the form $(a + \delta a_i)_i e^{j(\phi + \delta\phi_i - \frac{\pi}{2})} \hat{w}(\nu - \nu_i + \delta\nu_i)$ are subtracted from the FT of equation (1), so that one cannot hope the sum in equation (1) to be perfectly dismantled. The algorithm is biased. Moreover, if we continue iterating in Fig. 2, an artifact peak will be

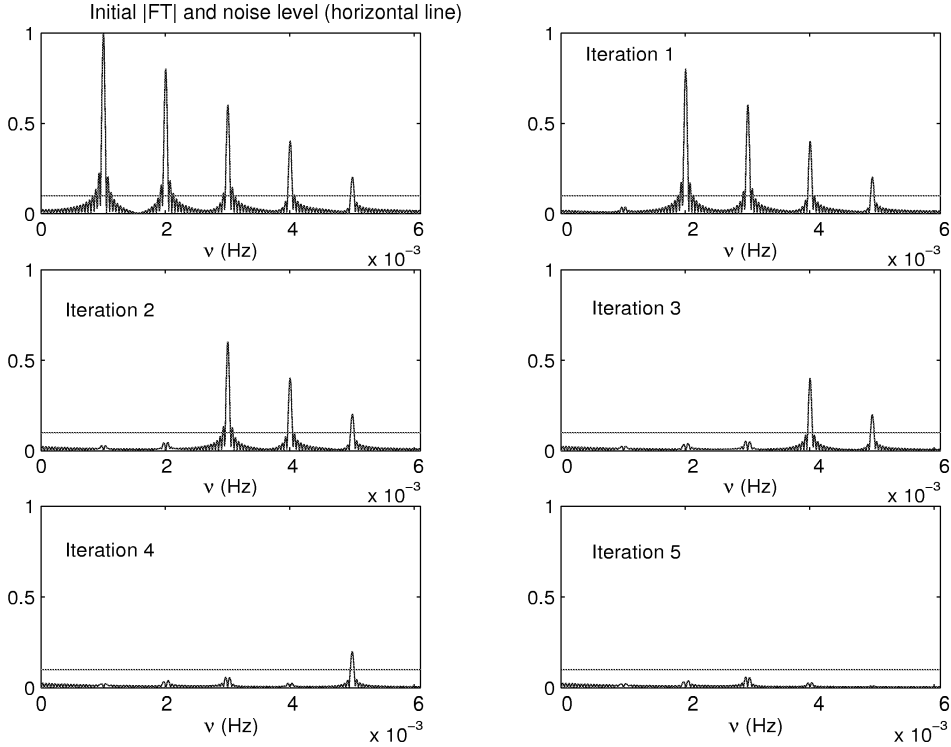


Figure 2. CLEAN's results for a 5 sinusoids (frequencies at 1, 2, 3, 4 and 5 mHz) signal.

found somewhere around 3 mHz, and then further ones if we continue iterating. For real signals exhibiting closely separated modes and/or gaps and/or noise, CLEAN's drawbacks are emphasized. Foster (1995) proposed an improved version of CLEAN (CLEANEST) which suffers however from the same impairments when there are gaps (Janot-Pacheco *et al.* 1999). Note also that a proper noise level estimation is capital. Many techniques exist for that purpose, see e.g., Roques *et al.* (1999) and references therein. In the particular case of WET data, Castanheira *et al.* (2004) use the fact that FT peaks whose amplitude is above four times the square root of the average power has a 1 over 1000 probability to be noise. Once high amplitude peaks have been removed, the noise in the residual FT can be estimated as the square root of the average power (Kepler 1993). For WET data, it is then useful to weight differently several runs: the higher the signal to noise ratio of the run, the larger the corresponding weight (Handler *et al.* 2002).

3.2 Several Peaks Deconvolution (SPD)

Since the reciprocal peaks' interferences contribute to CLEAN's bias, one may seek to estimate the set of $\{a_i, \phi\}$ jointly. This was proposed by Pfeiffer (1993). The basic idea is as follows:

- (1) Determine a number of K main modes and associated frequencies $\{\tilde{\nu}_i\}$,
- (2) Compute *jointly* the $\{\tilde{a}_i, \tilde{\phi}_i\}$ ("one-shot" (instead of iterative) approach),
- (3) Repeat 1 and 2 for other sets of $\{\tilde{\nu}_i\}$ to find the best fit w.r.t. to the data.

The signal model is unchanged, and the FT of the data is again modeled as in equation (1), in which only positive frequencies are considered (say N frequency points). For every frequency $\nu = 0, \dots, \nu_{\max}$, $\hat{s}(\nu)$ equals the sum of K contributions. The goal is to find the $\{z_i\} = \frac{a_i}{2} e^{j(\phi_i - \frac{\pi}{2})}$ which resolve this system of equations. The amplitudes and phases can be obtained by $a_i = 2|z_i|$ and $\phi_i = \arg\{z_i\} + \frac{\pi}{2}$. This problem can easily be put in matrix form: $\hat{W}Z = \hat{S}$, where $Z = [z_1 \ z_2 \ \dots \ z_K]^T$, $\hat{S} = [\hat{s}(0) \ \dots \ \hat{s}(\nu_{\max})]^T$, and $\hat{W}$ is the $(N \times K)$ *observation matrix*. The system being overdetermined (more data points N than peaks K), one may obtain an approximated solution by LS. The formal solution Z_0 is

$$Z_0 = \arg \min_Z \|\hat{S} - \hat{W}Z\|^2 \Rightarrow Z_0 = (\hat{W}\hat{W}^H)^{-1} \hat{W}^H \hat{S}, \quad (3)$$

where ^{H} denotes Hermitian transposition. As for CLEAN, the best LS fit of SPD may not lead to good solutions for noisy signals. Why?

4. The point of view of ill-posed problems

Usually, the determination of the oscillations' parameters is made by LS¹ fitting, either in the frequency domain (see above), or in the temporal domain² (prewhitening, close to CLEAN, see Ponman 1981). In both cases, the problem is equivalent to equation (3). The solution minimizes the MSE, but how reliable is this solution? In other words, how close are the estimated parameters from the actual ones?

That an LS fit (or any other data fitting technique) may not be meaningful is something that any astronomer is probably aware of³, and the precautions regarding the use of LS have been discussed elsewhere (see Isobe *et al.* (1990) for applications in Astronomy). It is maybe worth illustrating differently the meaning and implications of the questions above. Let us take for that purpose the simple example of a 2×2 observation matrix $\hat{W}$, and a 2 points data vector $\hat{S}$ for the system $\hat{W}Z = \hat{S}$. In this case the solution Z_0 is the intersection of two straight lines, see Fig. 3.

The sensitivity to noise can be more generally formulated as follows. Denote by λ_i (resp. v_i) the singular values (resp. vectors) of $\hat{W}$, then the LS solution of equation (3) becomes $Z_0 = \sum_{i=1}^K \frac{1}{\lambda_i^2} (v_i^T \hat{W}^T \hat{S}) v_i$. We can always write the data vector as $\hat{S} = \hat{S}_{\text{star}} + \hat{S}_{\text{noise}}$, where $\hat{S}_{\text{star}}$ is the signal delivered by the star and $\hat{S}_{\text{noise}}$ is the contribution of all the possible noise sources of section 2.2. Then we have

$$Z_0 = \sum_{i=1}^K \frac{1}{\lambda_i^2} (v_i^T \hat{W}^T \hat{S}_{\text{star}}) v_i + \sum_{i=1}^K \frac{1}{\lambda_i^2} (v_i^T \hat{W}^T \hat{S}_{\text{noise}}) v_i. \quad (4)$$

The second term in (4) shows that for small λ_i , the reconstructed LS solution suffers from *noise amplification*. Furthermore, denoting by δX a perturbation on a vector X

¹Invented by Gauss and Laplace in the beginning of the XIXth century for astronomical applications.

²Combination of linear and non-linear LS are sometimes used, see e.g., Kurz *et al.* (1997) for a detailed example.

³A major justification of LS is to yield the same estimate as the Maximum Likelihood Estimate for Gaussian residuals — a condition not often met in practice.

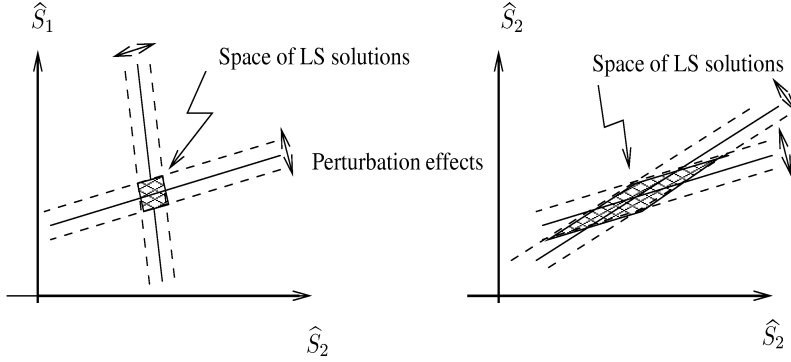


Figure 3. Illustration of the instability phenomenon in ill-conditioned problems. Left: The LS solution Z_0 is at the intersection of the two solid lines. If some perturbations occur on this setting (noise in the data), the resulting LS solutions will be limited to the dashed zone. Right: $\hat{W}$ has been changed. The LS solution Z_0 is still at the intersection of the two solid lines. With the same perturbation power as in the left panel, the corresponding LS solutions are very unstable (S. Jankov 2003, personal communication).

and by $\|X\|$ its norm, one can show that $(\delta Z_0)/\|Z_0\| \leq (\lambda_{\max}/\lambda_{\min})(\delta \hat{S}/\|\hat{S}\|)$. The ratio $\lambda_{\max}/\lambda_{\min}$ is usually referred to as the *condition number* of the matrix $\hat{W}$. The former inequality shows that little perturbations on the data create space for strongly different LS solutions. In asteroseismology, one obtains typically good conditioning numbers for one (clear) night ($\lambda_{\max}/\lambda_{\min} \approx 1$ or 2), whereas this number becomes very large for two nights ($\lambda_{\max}/\lambda_{\min} \approx$ a few hundreds). The effects of gaps are to introduce zeros in the $\hat{W}$ matrix whose columns tend to become similar. The weakest singular values tend to be even weaker, leading to the noise amplification in (4) and to a situation comparable to that illustrated in Fig. 3, right panel. The solutions $\{v_i, a_i, \phi_i\}$ obtained in this case are not trustworthy: because of noise and gaps, the data are just too different from the model, which corresponds to good observation conditions. Note that the analysis above does not leave us with doubt only: we have some mean to measure in a precise, mathematical form, however far our results may be from the actual parameters.

This kind of problems are named *ill-posed* or *ill-conditioned* problems, which means that the solution to a proposed problem may not exist, may not be unique, and may not be stable. In this case, many remedies exist, whose roots lie in the analyses above. Improved LS-like methods have been on the one hand designed a long time ago: e.g., reweighted LS allowing one to weight the data points/sets according to some confidence criterion; total LS which accounts for the presence of noise both in the data and the observation matrix (Golub & Van Loan 1980); different periodograms may also be used, e.g., Lomb-Scargle periodogram (Scargle 1982); L_1 norm in equation (3) can be minimized (which increases the robustness); and combinations thereof (see Branham (1990) for a review). Starting with Tikhonov & Arsenin (1977) on the other hand, a series of signal restoration methods have been conceived. For example, one may truncate the singular value decomposition so that the reconstructed solution in equation (4) involves only sufficiently large eigenvalues; *a priori* constraints on the reconstructed solutions (e.g., no aliases allowed, or smooth spectra) can be imposed (*regularization* solutions in Bayesian approaches), etc. Such deconvolution problems constitute an active field of research in the signal processing community (see the IEEE

literature) and, to a less extent, in asteroseismology (see e.g., Roques *et al.* 1999). In many approaches, the deconvolution is often performed over restricted frequency supports only (*i.e.*, where signal is large w.r.t. noise) so that the conditioning is better. But one understands that in this case the question becomes: how to separate signal and noise subspaces? Indeed, we miss knowledge about the part of the spectrum in which the pulsations may occur, since this is precisely what we are looking for. An important branch of these developments is supported by time-frequency representations, the most widely used being Wavelets (Daubechies 1992) and Matching Pursuit (Mallat & Zhang 1993). In these representations the time information is not lost (as in Fourier spectra). This feature is particularly attractive for noise discrimination (see Donoho & Johnstone 1994) and amplitude modulation detection. Roques *et al.* (1999) have used MP to detect which frequency intervals contain signal, and to perform deconvolution on the corresponding frequency supports. Such approaches allow both a high resolution restoration of $m(t)$ – even better than FT on some frequency intervals – along with a precise evaluation of the reliability of the solution.

5. Conclusions

Because uncooperative weather, atmospheric and instrumental noises are unavoidable in astronomical observations, data gaps and other noise sources remain major limiting factors for the accuracy of frequencies determination in asteroseismology. As such, their effects deserve to be most precisely understood. When analyzing light curves, astronomers traditionally use LS techniques and FT representation. Being aware of CLEAN-like method's impairments, the most careful ones use them for guidance only. In practice, the FT is often dismantled by hand, and it is analyzed according to the available information about the observational conditions, to the astronomers' previous knowledge of the star and, last but not least, to their own experience. In the parlance of Bayes, this is named *a priori* information. Indeed, such traditional methods often work, otherwise asteroseismology would not have shown so many excellent results – those reported in this special issue for example. However, many other efficient signal analysis tools exist – and many remain to be developed. These are interesting, regarding at least three points.

Firstly, these tools can provide a precise evaluation of the confidence level of the results.

Secondly, they may be more appropriate than the traditional techniques discussed above in the case of particular asteroseismologic phenomena (e.g., wavelets/MP help analysing amplitude modulation and frequency drifts).

Thirdly, such approaches provide an effort towards an explicit, mathematical formulation of the *a priori* information that can be injected in the analysis — in contrast to the more implicit *savoir-faire* of experienced asteroseismologists.

References

- Balona, L. A. 2002, In *Observational Aspects of Pulsating B and A Stars*, ASP Conf. Series, **256**, 43.
- Branham, R. L. Jr., 1990, In *Errors, Bias and Uncertainties in Astronomy*, Cambridge Univ. Press, p. 171.
- Breger, M., Handler, G., Nather, R. E. 1995, *A & AS*, **297**, 473.
- Castanheira, B. G. *et al.* 2004, *A & A*, **413**, 623.

- Christensen-Dalsgaard, J., Gough, D. O. 1982, *MNRAS*, **198**, 141.
- Christensen-Dalsgaard, J. 2003, *Lecture Notes on Stellar Oscillations*, available at www.asteroseismology.org
- Daubechies, I. 1992, *Ten Lectures on Wavelets*, SIAM, Philadelphia.
- Donoho, D., Johnstone, I. M. 1994, *Biometrika*, **81**, 425.
- Dravins, D., Lennart, L., Mezey, E. 1998, *Publications of the ASP*, **110**, 610.
- Finley, D. S., Koester, D., Basri, G. 1997, *ApJ*, **488**, 375.
- Foster, G. 1995, *AJ*, **109**, 1889.
- Golub, G. H., Van Loan, C. F. 1980, *SIAM J. Numer. Anal.*, **17**, 883.
- Handler, G., Romero-Colmenero, E., Montgomery, M. H. 2002, *MNRAS*, **335**, 399.
- Handler, G. 2004, *Com. in Asteroseismology*, **145**, 71.
- Hogbom, J. A. 1974, *A & AS*, **15**, 417.
- Isobe, T., Feigelson, E. D., Akritas, M. G., Babu, G. J. 1990, In *Errors, Bias and Uncertainties in Astronomy*, Cambridge Univ. Press, p. 25.
- Janot-Pacheco, E., Jankov, S., Leister, N. V., Hubert, A. M., Floquet, M. 1999, *A & A*, **137**, 407.
- Kepler, S. O. 1993, *Baltic Astron.*, **2**, 515.
- Kurz, D. W., van Wyk, F., Roberts, G., Marang, F., Handler, G., Medupe, R., Kilkeny, D. 1997, *MNRAS*, **287**, 69.
- Ledoux, P. 1951, *ApJ*, **114**, 191.
- Martinez, P. 1993, In *The Cape Rapidly Oscillating Ap Star Survey*, Ph.D Th., Univ. of Cape Town.
- Michel, E. *et al.* 1992, *A & A*, **255**, 139.
- Nather, R., Winget, D., Clemens, J., Hansen, C., Hine, B. 1990, *ApJ*, **361**, 309.
- Mallat, S., Zhang, Z. 1993, *IEEE Trans. on Sig. Proc.*, **41**, 12, 33.
- Pfeiffer, B. 1993, *Balt. Astronomy*, **2**, 538.
- Ponman, T. 1981, *MNRAS*, **196**, 583.
- Poretti, E. 2002, In *Observational Aspects of Pulsating B and A Stars*, *ASP Conf. Series*, **256**, 269.
- Roques, S., Mary, D., Dolez, N. 1998, *A Half Century of Stellar Pulsation Interpretations*, *ASP Conf. Series*, 447.
- Roques, S., Serre, B., Dolez, N. 1999, *Mon. Not. R. Astron. Soc.*, **308**, 876.
- Samadi, R., Nordlund, A., Stein, R. F., Goupil, M. J., Roxburgh, I. 2003, *A&A*, **403**, 303.
- Scargle, J. D. 1982, *ApJ*, **263**, 835.
- Scwarz, U. J. 1978, *A & A*, **65**, 345.
- Tikhonov, A. N., Arsenin, V. Y. 1977, *Solutions of ill-posed problems*, Winston, Washington DC.
- Unno, W., Osaki, Y., Ando, H., Saio, H., Shibahashi, H. 1989, *Nonradial oscillations of stars (2nd edition)*, Univ. of Tokyo Press, Tokyo.
- Warner, B. 1988, *High-speed astron. photometry*, Cambridge Univ. Press, Cambridge.
- Young, A.T. 1967, *Astron. J.*, **72**, 747.

Collaborative Research of Open Star Clusters

Alisher S. Hojaev

*Ulugh Beg Astronomical Institute, Center for Space Research, Uzbek Academy of Sciences,
Tashkent, Uzbekistan.*

e-mail: ash@astrin.uzsci.net

Abstract. Preliminary results on observations of open clusters are presented. The project has been initiated in the framework of the Uzbek–Taiwan and Taiwan–Baltic collaboration, mainly to upgrade and make use of facilities at Maidanak Observatory. We present detailed, multi-wavelength studies of the young cluster NGC 6823 and the associated complex nebulosity, to diagnose the young stellar population and star formation history in the region. In addition, 7 compact open clusters have been monitored for stellar variability. We show how observations like these could feasibly be used to look for exoplanet transit events. We also expect to join the Whole-Earth Telescope effort in future campaigns for asteroseismology.

Key words. Galaxy: open clusters and associations, general—method: observational.

1. Introduction

Open star clusters are one of the most important elemental components in the hierarchic structure of the Galaxy and of most other galaxies as well (see, e.g., Efremov 1989 and references herein). The stars forming a cluster are mainly born conjointly from the densest gas clumps in molecular clouds and therefore should have similar distances, ages, chemical composition (metallicity) and kinematics. Distributing throughout and around the Galactic disk, open star clusters are useful tools to probe the Galactic structure, dynamics and star formation history. According to the latest estimations, about 1700 open star clusters and cluster candidates are known in our Galaxy (WEBDA, Ruprecht *et al.* 1981; Lauberts 1982; Lynga 1987; Dias *et al.* 2002, 2004) but most of them are studied poorly or have not even been observed at all. At present due to the invention of more efficient two-dimensional detectors best adapted for the observations of star clusters, CCD observing is becoming the preferred technique. It replaces both the photoelectric and photographic photometry. At sites with quite good seeing conditions but equipped with small size telescopes the multicolor photometric and time-series monitoring observations of compact open clusters seems more worth pursuing than their spectroscopic surveys.

2. Observations and data processing

2.1 NGC 6823

The observations of complex star cluster NGC 6823 with bright nebula NGC 6820 were started in 2001 at Mt. Maidanak (Uzbekistan) on the 1.5 m and one of the 0.6 m telescopes. The Zeiss 0.6 m telescope equipped with a high sensitive water-cooled CCD-camera, Apogee AP-8e (SITe SI-003AB 1024×1024) was used for mosaicing photometric survey in R, I and H_α ($\lambda = 656.3$ nm; $\delta\lambda = 30$ nm) bands of the entire region. The 1.5 m AZT-22 telescope equipped with LN-cooled CCD-camera, BroCam (SITe-005 2000×800) was used in UBVR and H_α for small regions that called for careful examination in the images taken by the 0.6 m telescope. The standard stars (PG 2213, SA-94, 109, 110, 114, 115, etc.) were observed for calibration each run, sometimes a few during the night.

Some spectra of the young star candidates with dispersion 50 and 200 Å/mm were obtained in 2001 and 2002 with the spectrometer on the 2.16 m telescope of Beijing Observatory in China. Besides the wide field narrow band photometry images of the entire region NGC6823/20 were obtained with Beijing 60/90 Schmidt telescope in 2002.

The data on these observations are given in Table 1. In addition to these data 5 new frames were obtained by Beijing 60/90 Schmidt telescope in red continuum around H_α in May 2004 and a few photometric images of NGC 6823 core were obtained with Lulin 1 m telescope in Taiwan in November 2004.

Table 1. NGC 6823/20 observation data.

Maidanak 0.6 m	Maidanak 1.5 m	Beijing 60/90 Schmidt	Beijing 2.16 m
Aug 12–18, 2001 146 frames RIH $_\alpha$	Aug 08–17, 2001 53 frames UBVRH $_\alpha$	Sep 05–07, 2002 41 frames t, i, o	Oct 27–30, 2002 9 frames spectra

Notes: t-band = H_α ; i-band = $\lambda 6660$ Å and o-band = $\lambda 9100$ Å.

2.2 Time-series monitoring

A sample has included the compact unstudied open clusters or cluster candidates extracted from Dias *et al.* (2002), which clearly show the concentration of stars to most probably be a real cluster. The observations were carried out with Zeiss 0.6 m telescope (Maidanak, Uzbekistan) equipped with CCD-camera Apogee AP-10 (Thompson Atmel THX7899 2048×2048) in the period from July 31st to August 23rd, 2003. We have patrolled (continuously monitored one target during all night to obtain more dense time-series) seven open clusters with the following time allocation for each: IC 4996 (5 nights), NGC 7801 (3), King 1 (4), King 13 (3), King 18 (4), King 20 (3), and Berkeley 55 (3). The R and I filters were used. The standard areas around SA 113-342 and 112-275 were observed for calibration.

2.3 BVI photometric imaging

These observations were made at Lulin Observatory (Taiwan) in 2003, November 08–12 in frame of Taiwan–Baltic–Uzbek collaboration and ongoing projects at National

Central University of Taiwan where the author worked as visiting scientist at that time. 1 m telescope with CCD-camera AP-8e described above was used with Johnson BVI filters. 20 poorly studied open star clusters were imaged for multi-wavelength photometry (King 1, King 16, NGC 1624, Berkeley 24, Stock 20, Skiff 1, NGC 1798, NGC 2262, Stock 21, Stock 3, NGC 1857, Berkeley 25, Dias 1, NGC 657, NGC 2126, Berkeley 75, Czernik 2, Czernik 6, NGC 2192 and Berkeley 27). We used the standards at SA – 92, 96, 98, 112, 114, RU149 to derive transformation parameters and to calibrate our photometric data.

2.4 Data reduction and calibration

The basic CCD-data reductions for the bias, dark and flat corrections have been done using the IRAF package routines for the NGC 6823, IC 4996, NGC 7801, King 1, King 13, King 18, King 20 and Berkeley 55 data. We have also made the proper WCS-registration of NGC 6823 frames using the scripts developed by Daisuke Kinoshita. Typical error box was not larger than $0.2''$ for each coordinate. For the extraction of sources, their analysis and photometry of stars we used the SE v2.2 by Emmanuel Bertin. The standard error of photometry (σ) in R and I bands is shown in Fig. 1.

In the case of NGC 6823 we applied all processing procedures to each frame of mosaic and analyzed every overlapped frame for consistency. The conformity of the magnitudes of stars in the overlapping areas is not larger than the standard error of photometry. This has allowed us to combine all separate catalogs of mosaic into one master catalog for each filter. The special script was written for comparison and cross-identification of catalogs. The resulting RI master catalog in the outcome consists of 3183 stars with measured R and R – I values.

3. Results, analysis and discussion

NGC 6823 is a sufficiently young, rich and compact open star cluster embedded into bright nebulosity with prominent dark trunks and globules. The core of the cluster as

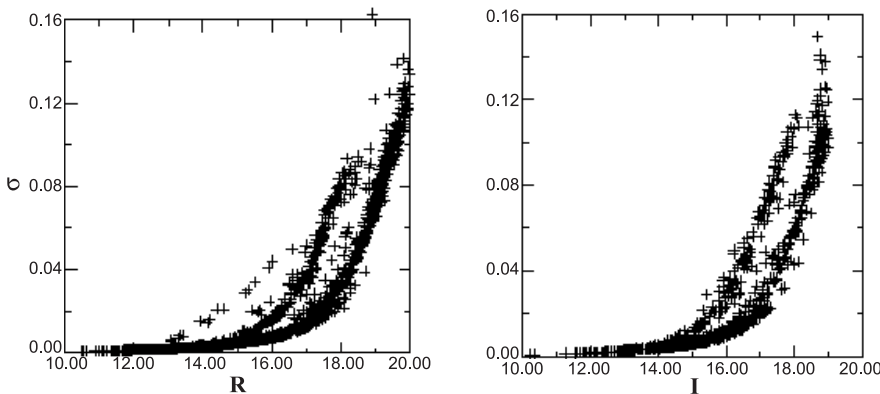


Figure 1. Standard error (σ) for R (left panel) and I (right panel) magnitudes.

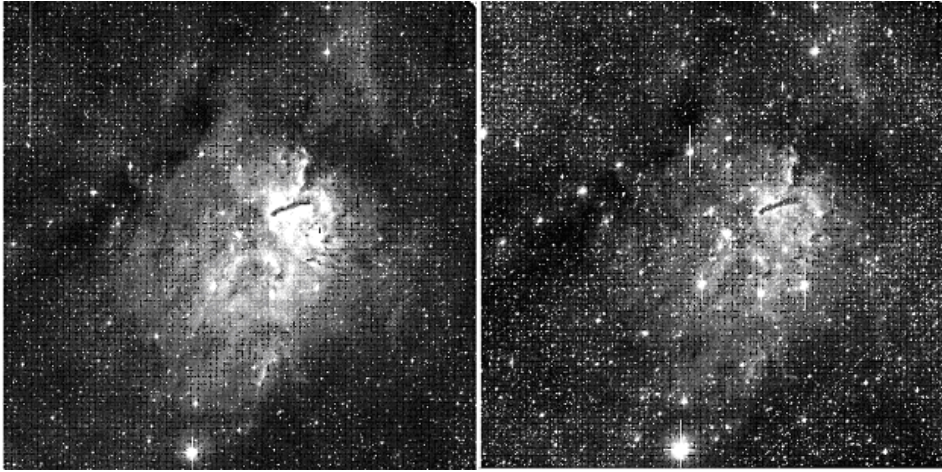


Figure 2. Left panel – the i-band ($\lambda = 6660 \text{ \AA}$) and right panel – the o-band ($\lambda = 9100 \text{ \AA}$) images of NGC 6823/6820 obtained by Beijing Observatory 60/90 Schmidt telescope. N – at the top; E – at the right. Open cluster NGC 6823 is located at the center of the images.

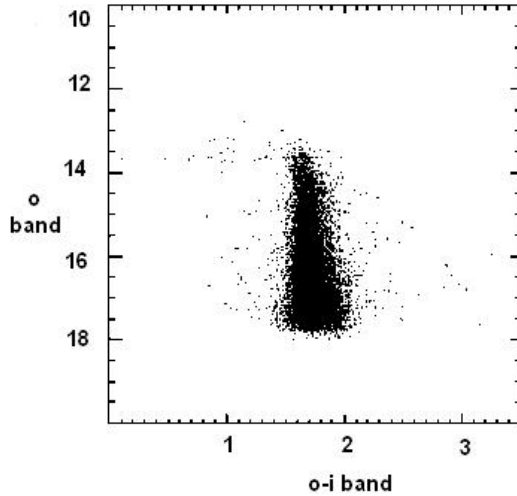
well as the nuclear part of giant OB stars association Vul OB1 connected with the cluster consists of a trapezium type multiple system. In common by most indicators this aggregate is very similar to the star formation complex in Orion. In Fig. 2, the narrow-band BATC image of star cluster NGC 6823 and bright HII nebula NGC 6820 is shown in o and i band passes; for more details on observational data, see Table 1. In both images the open cluster is located at the center of the picture.

Low and intermediate mass PMS stars (T Tauri stars of the late spectral type and early type Herbig emission-line stars, respectively) usually have a near-infrared (NIR) excess emission as compared to the normal stars of the same spectral types. Based on the 2MASS NIR colors and selection procedure outlined in Lee & Chen (2002) 60 young star candidates have been selected in the wide region around NGC 6823. The spectra were obtained for 9 most bright and suitable for spectral observations candidates (Table 1). Eight sources (*i.e.*, 89 per cent from the sample for which the spectral material was obtained) have been found to show clear H-alpha emission in their spectra, of which 5 are probable Herbig Ae/Be stars (HAEBES), and 3 are probable classic T Tauri stars (CTTS). The spectral energy distribution (SED) of these stars also has the shapes typical to the SED for corresponding classes of PMS stars. These young star candidates, and their 2MASS NIR magnitudes, are summarized in Table 2. These stars locate in the same parts of the entire region as known PMS stars and related objects as well as the cluster and OB association stars.

The o and i narrow band frames of NGC 6823/20 were used for photometric measurements. The instrumental magnitudes for stars were calibrated by PSF fitting. 12790 stars in i band image and 17903 stars in o band image have been extracted, and 11308 stars have both i band and o band data. The color-magnitude diagram (CMD) for o band magnitude versus o band minus i band color index is presented in Fig. 3. Only the stars with brightness fainter than 13 magnitudes were considered in o band. The bright (massive) stars in this area could not be measured for photometry because they are saturated. The presence of foreground and background stars in Fig. 3 makes the main-sequence wider.

Table 2. Some newly found young stars in the NGC 6823/6820 complex.

RA (J2000)	Dec (J2000)	Type	J	H	K _s
19 42 27.92	+23 05 14.7	CTTS	11.047	10.275	09.689
19 42 54.99	+23 24 14.8	CTTS	10.343	09.333	08.404
19 42 57.68	+23 22 52.1	CTTS	12.955	12.213	11.622
19 43 04.40	+23 18 48.7	HAeBeS	11.579	10.515	09.601
19 43 07.51	+23 26 02.9	HAeBeS	12.753	11.811	10.970
19 43 08.59	+23 25 45.7	HAeBeS	13.595	12.784	12.066
19 43 10.17	+23 25 31.1	HAeBeS	13.558	12.304	11.381
19 43 20.99	+23 19 02.3	HAeBeS	13.057	11.557	10.355

**Figure 3.** Color-magnitude diagram for o band and o band minus i band for stars in the region of NGC 6823/20.

According to recent estimations (see, for example, the Wilton Dias' latest statistics, Version 2.0 of June 2004, or WEBDA database) less than half of known open star clusters and cluster candidates have been studied at least photometrically to construct their CMD as well as color diagrams and derive the main parameters such as distances, reddening and ages. The kinematical or membership analysis based on radial velocities and proper motions of member stars has been made for about one tenth of the total number of open clusters. Chemical abundance ratio and [Fe/H] metallicity parameters were determined for even smaller numbers. The result of the statistical analysis is presented in Table 3.

Obviously this incompleteness points out that a large effort is still needed to improve the catalogue data. Moreover we reviewed the SIMBAD database and the literature for screening the search and study of variable stars or any kind of variability in open clusters. It turned out, that the number of such clusters is about a few dozens only and the variables in that clusters were as a rule, found accidentally. The few special

Table 3. Statistical analysis for open cluster parameters investigation.

Feature of clusters	Number	Fraction
Clusters with diameters	1688	0.999
Clusters with distances	740	0.438
Clusters with reddening	721	0.427
Clusters with ages	608	0.360
Clusters with distances, reddening and ages	594	0.351
Clusters with proper motions (PMs)	614	0.363
Clusters with radial velocities (RVs)	227	0.134
Clusters with PMs + RVs	217	0.128
Clusters with distances, ages, PMs and RVs	214	0.127
Clusters with abundances	131	0.078
Total	1690	1.000

monitoring runs were only reported on the search for variables in clusters other than the Pleiades, Coma Berencis, in Orion and so forth where the flare stars were looked for. At the same time the revealing of some kin variables in clusters will allow us to get more specific information about the cluster itself. For instance, the eclipsing binaries or cepheids make more exact the distance determinations, the discovery of PMS and flare stars – the evolutionary status and age. Consequently the monitoring of selected open clusters for variability is a worthwhile task. As a by-product of such time-series analysis the discovery of exoplanet transits might follow.

Acknowledgements

This work has been supported by STC Uzbekistan (grant F-2-2-5) and by grants of NSC Taiwan through the Taiwan-Baltic Foundation. The author is grateful to W. P. Chen, Zhou Xu, H. T. Lee, Z. W. Zhang, D. G. Semakov for the participation in observations and discussions. This research has made use of the SIMBAD database, operated at CDS, Strasbourg, France. The publication also makes use of data products from the Two Micron All Sky Survey (2MASS), which is a joint project of the University of Massachusetts and the Infrared Processing and Analysis Center/California Institute of Technology, funded by the National Aeronautics and Space Administration and the National Science Foundation. We also use data from Digitized Sky Survey which were produced at the Space Telescope Science Institute under U.S. Government grant NAG W-2166. The author is thankful to an anonymous referee for useful comments.

References

- Dias, W. S., Alessi, B. S., Moitinho, A., Lepine, J. R. D. 2002, *Astron. Astrophys.*, **389**, 871.
Dias, W. S., Alessi, B. S., Moitinho, A., Lepine, J. R. D. 2004, *Computer Based New Catalogue of Open Cluster Data*, Strasbourg: CDS.
Efremov, Yu. N. 1989, *Sites of Star Formation in Galaxies: Star Complexes and Spiral Arms*, Moscow, Nauka Publ. House, p. 246.
Lauberts, A. 1982, *ESO/Uppsala survey of the ESO(B) atlas*, Garching: European Southern Observatory.

- Lee, H. T., Chen, W. P. 2002, *Proceed. 8th IAU Asian-Pacific Regional Meeting*, Tokyo, 101.
- Lynga, G. 1987, *Computer Based Catalogue of Open Cluster Data*, 5th edn, Lund Observatory (Strasbourg: CDS).
- Ruprecht, J., Balazs, B., White, R. E. 1981, *Catalogue of Open Clusters and Associations*, Akademiai Kiado, Publ. House Hungarian Acad. Sciences, Budapest.

Results from the Whole Earth Telescope – Indian Contributions

S. Seetha^{1*}, B. N. Ashoka¹ & T. M. K. Marar²

¹*Space Astronomy and Instrumentation Division, ISRO Satellite Centre, Airport Road, Bangalore 560 017, India.*

²*'Swathi' near Valiyalukkal Temple, Kanimangalam, Thrissur 680 027, India.*

**e-mail: seetha@isac.ernet.in*

Abstract. The Indian team at ISRO has been part of the Whole Earth Telescope (WET) team since 1988 when we first participated in the campaign on V471 Tau. We have been part of many other observing campaigns over the last decade. This presentation traces the circumstances leading to our joining the WET team and how useful the coverage from the Indian longitude has been. The results of several pulsators from the WET runs during which we participated are also described. These include PG1159-035 the prototype of the GW Vir type of stars, RE J 0751+14 a cataclysmic variable, PG 1336-018 a binary with an sdB pulsator and finally HR 1217 a roAp star. The paper concludes with what the limitations are in our observations and how we can overcome them in the future.

Key words. Stars: pulsating—stars: individual: PG1159-035—stars: individual: RE J 0751+14—stars: individual: PG 1336-018—stars: individual: HR 1217.

1. Introduction

The Whole Earth Telescope (WET) is a world wide network of astronomers whose main goal is to study pulsating stars and to resolve the pulsations at the best ever frequency resolution. In order to achieve this purpose, the members of this collaboration try to obtain uninterrupted data with an improved signal to noise ratio on specific target stars. This leads to the detection of multiple frequencies to very low amplitude levels (a few millimag going down to a few micromag) and the resolving of closely spaced (to few μ Hertz) frequencies. The data set appears as though the observations on the star was a continuation from one observing site to another as the star moved westwards and as though the whole Earth served as an observing site — hence the name. The concept of WET was mooted in 1986 and originated in the precincts of the University of Texas at Austin, USA. In 1995, it was affiliated to the IITAP – International Institute of Theoretical Physics – at Iowa State University. The WET is headed by a Director who is assisted in all decision matters by the Council Of Wise (COW) (see <http://wet.physics.iastate.edu>). However, it is a fairly democratic body with provision for all collaborators to interact and have their say. The first observational campaign was conducted in 1988 on PG 1346+082 with six observatories participating. Over the years there have been astronomers from 20 observatories who have participated

in various campaigns and 24 campaigns have so far been conducted. The WET campaigns are denoted by XCOV n where ' n ' is the number of the campaign. The Asian longitudes are very crucial because almost half of this area in terms of latitude is covered by the Indian Ocean and in the remaining land area there are not sufficient numbers of observatories with the necessary observing conditions. Hence, the observational coverage from the Asian longitudes are exigent to fill critical gaps in the data sets. India joined the WET organisation during the second campaign in November 1988 (XCOV2) to observe V 471 Tau. Since then we have contributed data sets during 7 campaigns.

2. Indian participation

Many astronomers in India have been involved in the observation of variable stars at various observatories. Members of our division (SAID; earlier known as Technical Physics Division) at ISRO Satellite Centre were involved in fast photometry of cataclysmic variables. Well aware of the problems connected with single channel photometers, we decided to build a two-star photometer. Our preliminary design was made in 1984. During the IAU general assembly held at New Delhi, India, in 1985, we got the chance to meet other astronomers who had built or were planning to build two/three channel photometers. Some of them are the founder members of WET. Having realised the two-star photometer in 1987 (Venkat Rao *et al.* 1990; also see Ashoka *et al.* 2001) we were in a position to test it in 1988. With the provision of the interface card and the Quilt-9 software (Nather *et al.* 1990) from our WET colleagues and the availability of PCs around the same time, everything fell into place for our joining the WET campaigns. Variable star photometry became fun with the capability to observe the light curve in real time. Since then we have participated in over 7 campaigns and have used telescopes at Vainu Bappu Observatory at Kavalur in South India (lat = $12^{\circ}34'.5$, long = $78^{\circ}49'.5$) as well as at ARIES (lat = $29^{\circ}, 22'$, long = $79^{\circ}, 27'$; erstwhile State Observatory Nainital).

3. Asteroseismology with the WET

The main goal of the WET has been asteroseismological studies of pulsating variables. Many of these variables pulsate in non-radial modes exhibiting multiple, closely-spaced frequencies. The aim is to resolve these frequencies and evolve a theoretical model which simulates the observed frequencies. Essentially, the observationally resolved frequencies can help to fine tune theoretical models to fit observations so that various physical parameters of these pulsating stars can be estimated. For most classes of variables, the same estimates can be extended to non-variables also as a class. The physical parameters which can thus be estimated using asteroseismological tools and which have been best demonstrated for the ' g ' modes exhibited in white dwarf pulsators are as follows:

- The rotation rate of the star: by the ' m ' splitting or rotational modulation of pulsations.
- The magnetic field strength: by the asymmetry in ' m ' spacing.
- Differential rotation rate: change in ' m ' splitting for different modes (different locations in the star).

- The stellar mass: from the mean period spacing of same ‘ l ’ but different ‘ k ’.
- Core composition by the mean period spacing of different adjacent ‘ k ’ values.
- Surface layer masses: by deviations from the mean period spacing.

Although the maximum success rate of applying asteroseismological methods has been achieved with white dwarfs, several other pulsation studies have been possible with WET used to observe other classes of pulsating variables also. We will describe in this paper a few of the results from WET on different classes of objects, selecting stars which were observed from the Indian longitudes. This however is NOT a summary of the overall results of the WET. We choose for the purpose of this paper, four targets:

- PG 1159-035 – a DOV white dwarf; first target for which most of the results of asteroseismology were demonstrated.
- PG 1336-018 – an eclipsing sdBV system where the sensitivity limits of the WET enabled the detection of very low amplitude pulsations even during eclipse.
- REJ 0751+14 – a cataclysmic variable, where long duration allowed for observing changes in the pulse profile.
- HR 1217 – a roAp star, where a theoretically predicted frequency was searched for and found.

3.1 PG 1159-035

PG1159-035 also known as GW Vir, or McGraw’s star was discovered to be a variable in 1979 (McGraw *et al.* 1979). It exhibits a predominant period at 516s. Further co-ordinated observations demonstrated that although the 516s is the predominant frequency, there are several groups of distinct frequencies clearly indicating that the star is multiperiodic. (Winget *et al.* 1985; Winget 1988). Based on these observations it was proposed that the star is pulsating in the ‘ $l = 1$ ’ or ‘ $l = 3$ ’ mode for a 0.6 solar mass white dwarf (Kawaler 1988). The change in period of the 516s or dp/dt was estimated to be $(-2.4 \pm 0.4) \times 10^{-11} \text{ss}^{-1}$ (Winget and Kepler 1988).

Knowing well that this star is multiperiodic, a Whole Earth Telescope campaign was conducted in March 1988 (XCOV3) in which 9 observatories were involved and the results were published by Winget *et al.* (1991). It was proven with this campaign that the star is indeed a rich multimode pulsator. Over 121 peaks were evident in the Fourier transform of which 101 modes were identified (Winget *et al.* 1991). Several ‘ $l = 1$ ’ modes and a few ‘ $l = 2$ ’ modes were identified. The fine splitting for both these sets indicated that the rotation period of the star is $1.38 \pm 0.01d$.

The non-uniformity in the splitting, which was less than 5%, was attributed to a magnetic field strength of less than 6000G. The determination of the mean period spacing for modes with the same ‘ l ’ but different values of ‘ k ’, enabled us to estimate the mass to be 0.586 ± 0.003 solar mass.

It was also concluded that the atmosphere of PG 1159-035 is compositionally stratified with a very thin layer of helium and even a thinner layer of hydrogen. Although the 516s period was identified to be an ‘ $m = +1$ ’ mode the dp/dt derived from earlier data was consistent with the WET observations and the value of dp/dt was refined to be $(-2.49 \pm 0.06) \times 10^{-11} \text{ss}^{-1}$; PG1159-035 has been observed several times after this campaign, using WET in 1993, 2000 and 2002. The frequencies of the resolved modes appear stable, although the amplitudes do vary, as is the case with several pulsators. However, all issues of PG1159-035 are not completely solved. Evolutionary

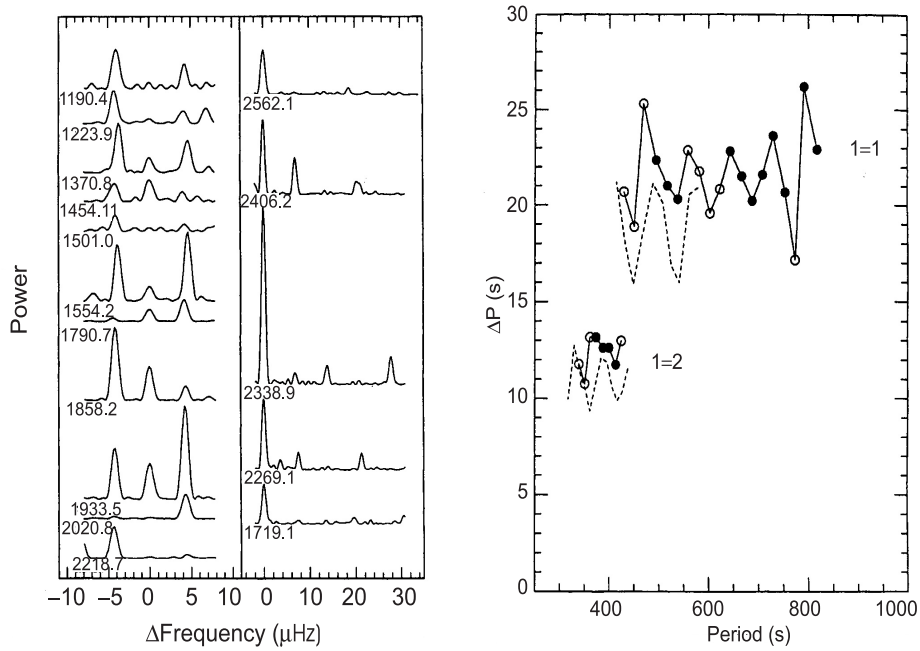


Figure 1. Figure on the left shows the $l = 1$ and $l = 2$ splittings of different modes stacked one above the other for viewing convenience. The figure on the right shows the same modes on period scales to show the approximately equal period spacing of the g modes (from Winget *et al.* 1991).

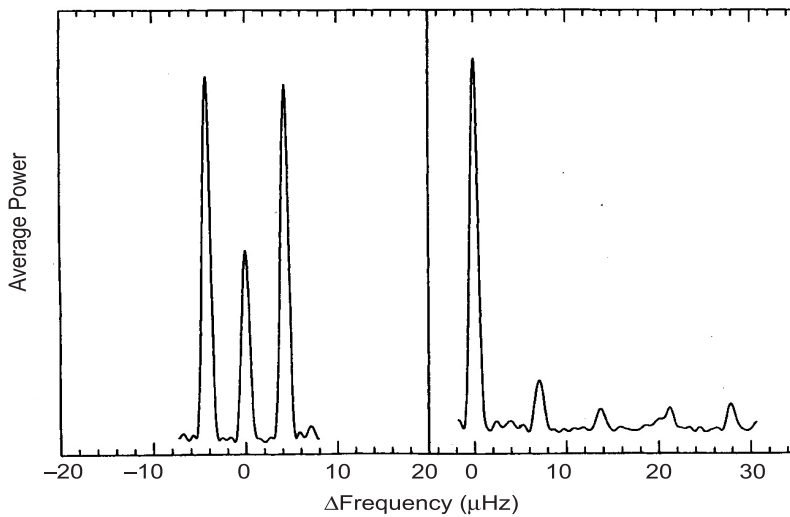


Figure 2. The average of all the $l = 1$ and $l = 2$ modes (from Winget *et al.* 1991).

models predict that dp/dt would be negative for trapped modes like the 516s and the 539s modes but would be positive for non-trapped like the 495s and 451s modes (Kawaler and Bradley 1994). Later direct measures (Costa and Kepler 1997) of phase of 516s period indicate that dp/dt for the 516s period itself could be positive and very

large $((+13 \pm 2.6) \times 10^{-11} \text{ss}^{-1})$. The evolution of the periods with time is an issue which will need many more observations before it is settled.

3.2 PG 1336-018

Also known as NY Vir, this system is a close eclipsing binary with a binary period of 2.4 hr (0.101d). One of the components of the binary is a pulsating B subdwarf (sdBV). These are also known as EC 14026 variables after the first prototype discovered in this group (Kilkenny *et al.* 1997). Pulsations were discovered in PG1336-018 with two frequencies at 5.4 mHz (184s) and 7.1 mHz (141s) being predominant (Kilkenny *et al.* 1998). Amplitude modulation and bands of power at other frequencies indicated that this star could be multiperiodic. PG 1336-018 was observed in April 1999 (XCOV17) with the WET and 172 hours of observation with 47% coverage was obtained. More than 28 frequencies were detected with 20 frequencies being identified in the frequency range of 4000 to 8000 μHz down to an amplitude of 0.003 modulation intensity (mi) level (Kilkenny *et al.* 2003). The predominant frequency around 5400 μHz itself was resolved into several frequencies. Frequency structure changes were observed on time scales as short as one day. This may be caused by unresolved modes or amplitude changes. Rotational effects were searched for and may have been present in two of the frequencies. Although 20 frequencies were identified, they could not be classified as specific modes due to the inadequacy in modeling. One of the aims of the WET campaign was to look for pulsations during eclipse and if well resolved, mode identification would be possible depending on the gradual decrease/increase in amplitude of the frequencies during ingress/egress.

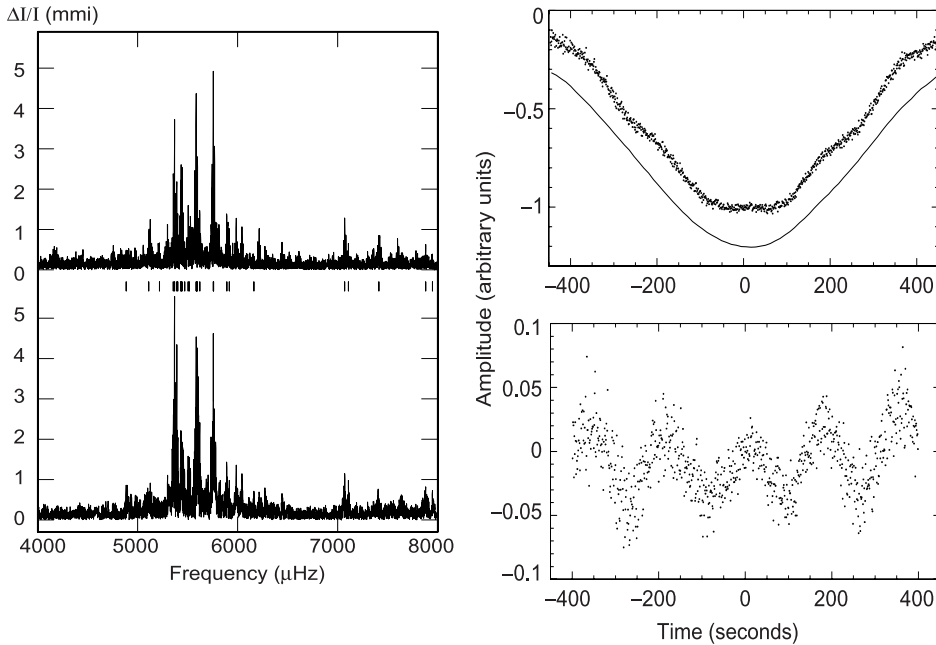


Figure 3. The figure on the left shows the frequency spectrum of PG 1336-018 obtained during two halves of the WET run in April 1999. The figure on the right shows the presence of pulsations during the eclipse after removing the eclipse profile using a fit (from Kilkenny *et al.* 2003).

The inclination of the system is estimated to be 81 degrees, and therefore even during primary eclipse, pulsations were expected. One important result obtained from the WET campaign was that during the eclipse, pulsations were indeed observed. However, they could not be resolved into modes due to insufficient coverage of eclipses during the campaign.

3.3 RE J0751+14

Also known as PQ Gem, this is an intermediate polar and has been extensively observed in both optical and x-ray bands (Mason *et al.* 1992; Duck *et al.* 1994; Rosen *et al.* 1993; Hellier *et al.* 1994). The rotation period of the white dwarf in this binary system was determined to be 13.9 min with its harmonic also at 6.9 min observed in R and I bands. The presence of a beat period was evident but the period of the beat was observed to be 14.5 min or 15.2 min. The polarisation studies in this system indicated that it is more like a polar with the estimated magnetic field to be as high as 8–18 MG using cyclotron models (Piirola *et al.* 1993; see also Vath *et al.* 1996; Potter *et al.* 1997). This object was observed with the WET during February 1996 (XCOV13). The spin period and the synodic period were clearly identified at 833.39s (13.9 min; frequency = 1.199 mHz) and 872.25s (14.5 min; frequency = 1.146 mHz) (Marar *et al.* 1998). The harmonics of these two periods were also observed. Both these periods were stable during the WET run. Based on this, the orbital period was estimated to be 5.19 h.

The spin profile and the beat pulse profile as a function of the orbital period was calculated which showed changes in the accretion geometry. The object was studied extensively almost simultaneously by many of the authors mentioned above. The white dwarf in the system was found to be spinning down at a rate of $(+1.1 \times 10^{-10} \text{ss}^{-1})$ by Mason (1997), which was confirmed in the WET data set ($dp/dt = 9.2 \times 10^{-11} \text{ss}^{-1}$).

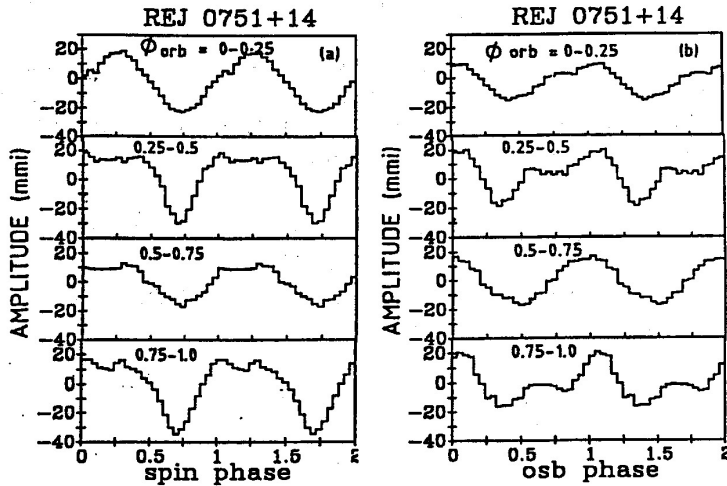


Figure 4. The variation in the profiles of the spin and orbital side band modulation as a function of orbital period for REJ 0751+14 (from Marar *et al.* 1998).

The object was studied by Potter *et al.* (1997) and extensively by spectroscopic methods by Hellier (1997). The spectroscopic results indicate that the system is indeed disc fed.

3.4 HR 1217

HR 1217 is a well studied rapidly oscillating Ap (roAp) star which exhibited six oscillation frequencies each of these showing rotational modulation (Kurtz *et al.* 1989). The spacing between the adjacent modes was found to be $34 \mu\text{Hz}$. This could be explained either as alternate even and odd modes which would imply a real splitting of $68 \mu\text{Hz}$ for the same values of ' l '; or the adjacent modes could be different radial orders (' n ' values) of the same ' l ' which would imply a real splitting of $34 \mu\text{Hz}$. This was resolved (Mathews *et al.* 1999) using the Hipparcos data and the asteroseismic and the Hipparcos luminosity matched only for alternate odd/even modes implying a real splitting of $68 \mu\text{Hz}$. This result however showed that whereas 5 adjacent frequency groups in the earlier data of the star could be explained by this splitting, the 6th frequency was at a frequency of only $3/4$ of this splitting frequency.

Since roAp stars have magnetic fields of several kilo Gauss, it was proposed that magneto-acoustic effects could be important and magnetic perturbation could result in variation in spacing (Cunha and Gough 2000; Bigot *et al.* 2000). Cunha (2001) predicted there would be an additional frequency with a splitting of around $34 \mu\text{Hz}$. This frequency was searched for during the WET run held in November 2000 (XCOV20).

An additional peak was detected at $2788.95 \mu\text{Hz}$, near the predicted frequency; this peak had an amplitude of 0.1 mma (Kurtz *et al.* 2002). It may be noted that this amplitude was well above the noise levels in the previous runs of 1986 and why it was not detected is still an open question. The amplitudes of various modes do change with time and this could be the reason for the non-detection of the new frequency in the 1986 dataset and its detection in 2000, which will be confirmed only with future observations. Meanwhile the MOST satellite has also observed this star and found almost the same frequency spectrum (see papers by Cunha; Kurtz; these proceedings and Kurtz *et al.* 2005).

4. Future of WET

It is clear that in the future we will need larger telescopes and will need to access fainter targets. At present, discussion groups are organised within WET to discuss the future of WET. This includes:

- Science issues which can be addressed by WET in the context of upcoming satellite experiments to observe pulsating variables; identification of niche problem areas in science which WET can address best;
- Funding issues as to how the WET should be operated in the future;
- Instrumentation issues as to whether all WET colleagues can be provided with identical instrumentation;
- Accessibility to fainter targets.

In particular the coverage of Asian longitudes has to be addressed, as this is becoming crucial from the point of view of gaps in data. The months of April to August are clouded over due to the monsoon in most parts of India. Alternate high altitude sites where weather patterns are different may prove very useful in this. For this purpose,

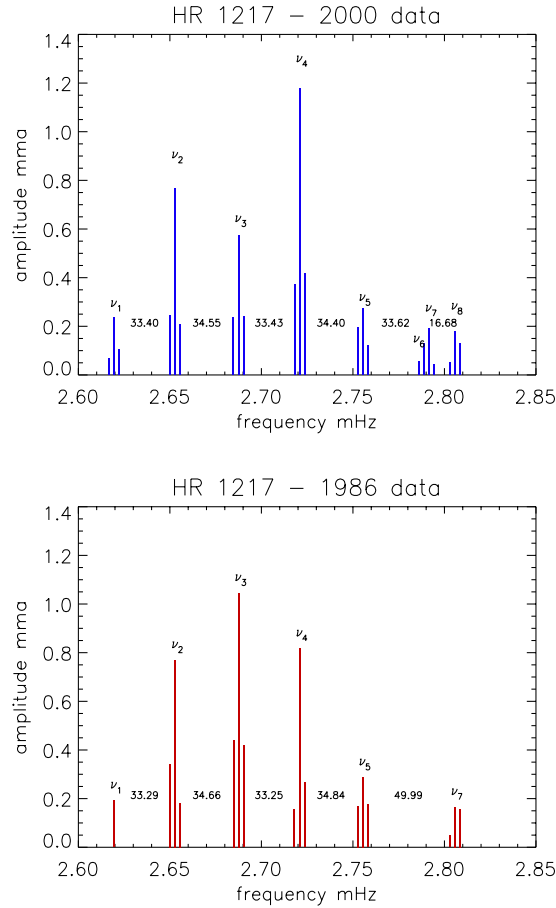


Figure 5. The difference in the Fourier spectra of the WET data set (top) and the 1986 dataset (bottom) (from Kurtz *et al.* 2002).

if bigger telescopes in better sites in India can be equipped with photometric specific instrumentation with good timekeeping (which is essential for time series photometry) and excellent tracking (since these objects are observed continuously for the whole night with as few breaks as possible), then these telescopes could all be included for future WET campaigns.

Acknowledgement

This paper was made possible by the contributions of all the WET team members.

References

- Ashoka, B. N. *et al.* 2001, *JapA*, **22**, 131.
 Bigot *et al.* 2000, *A&A*, **356**, 218.
 Costa, J. E. S., Kepler, S. O. 1997, *Baltic Astron.*, **7**, 83.

- Cunha, M. S. 2001, *MNRAS*, **325**, 373.
Cunha, M. S, this volume.
Cunha, M. S., Gough 2000, *MNRAS*, **319**, 1020.
Duck, S. R. *et al.* 1994, *MNRAS*, **271**, 372.
Hellier, C. 1997, *MNRAS*, **288**, 817.
Hellier, C., Ramseyer, T. F., Jablonski, F. J. 1994, *MNRAS*, **271**, L25.
Kawaler, S. D. 1988, In: *IAU Symp 123, Advances in Helio and Asteroseismology*, Dordrecht: Reidel, p. 329.
Kawaler, S. D., Bradley, P. A. 1994, *AJ*, **427**, 415.
Kilkenny D. *et al.* 1997, *MNRAS*, **285**, 640.
Kilkenny, D. *et al.* 1998, *MNRAS*, **296**, 329.
Kilkenny, D. *et al.* 2003, *MNRAS*, **345**, 834.
Kurtz, D. W. *et al.* 1989, *MNRAS*, **240**, 881.
Kurtz, D. W. *et al.* 2002, *MNRAS*, **330**, 57.
Kurtz, D. W. *et al.* 2005, *MNRAS*, **358**, 651.
Marar, T. M. K. *et al.* 1998, *Baltic Astr.*, **7**, 287.
Mason, K. O. 1997, *MNRAS*, **285**, 493.
Mason, K. O. *et al.* 1992, *MNRAS*, **258**, 749.
Matthews, J. M., Kurtz, D. W., Martinez, P. 1999, *ApJ*, **511**, 422.
Mcgraw, J. T. *et al.* 1979, In: *IAU colloq. 53, White dwarfs and variable degenerate stars*, Rochester N.Y.: University of Rochester Press, p. 377.
Nather, R. E. *et al.* 1990, *Astrophys. J.*, **361**, 309.
Pirola, V., Hakala, P., Coyne, G. V. 1993, *ApJ*, **410**, L107.
Potter, S. B. *et al.* 1997, *MNRAS*, **285**, 82.
Rosen, S. R., Mittaz, J. P. D., Halala, P. J. 1993, *MNRAS*, **264**, 171.
Vath, H., Chanmugam, G., Frank, J. 1996, *Astrophys. J.*, **457**, 407.
Venkat Rao, G. *et al.* 1990, *BASI*, **18**, 79.
Winget, D. E. 1988, In: *IAU Symp 123, Advances in Helio and Asteroseismology*, Dordrecht: Reidel, p. 305.
Winget D. E., Kepler, S. O. 1988, In: *Multimode stellar pulsations* (Budapest: Konkoly Observatory), p. 205.
Winget, D. E. *et al.* 1985, *Astrophys. J.*, **292**, 606.
Winget, D. E. *et al.* 1991, *Astrophys. J.*, **378**, 326.

Upgradation of MOT and its Relevance to WET Campaigns

Alisher S. Hojaev

*Ulugh Beg Astronomical Institute, Center for Space Research, Uzbek Academy of Sciences,
Tashkent, Uzbekistan.*

e-mail: ash@astrin.uzsci.net

Abstract. The Maidanak One-meter Telescope (MOT) is a Ritchey-Chretien telescope by Carl Zeiss, located atop Maidanak in Uzbekistan, where site parameters indicate excellent atmospheric seeing conditions. An effort to computerize the MOT, jointly made by an Uzbek, Taiwanese and Baltic consortium, is expected to complete in 2005. Monitoring the variability of star clusters will be among the first scientific projects to be carried out. Equipped with sensitive CCD cameras, the MOT, with its middle-Asia geographic location, will be a desirable addition to the Whole Earth Telescope (WET) network. We describe the upgrade engineering and instrumentation of the telescope and how the system can be used in future campaigns.

Key words. Telescopes—method: observational—stars: oscillation.

1. Introduction

Asteroseismology belongs to one of the main subjects of the Whole Earth Telescope (WET) activity. The principal aim of WET's Extended Coverage (XCov) is uninterrupted observations of the targets. For this purpose the observing sites ideally should be widely distributed in longitude. In the earlier years of the project such a longitudinal coverage was lacking. In an attempt to fill these gaps in the Eastern Hemisphere, Prof. Jan-Erik Solheim and Dr. Edmundas Meištas arranged a special exploration for potential sites in Central Asia in 1992, including possibility of Tien-Shan (Kazakhstan), Assy-Turgen (Kazakhstan), Sanglok (Tadjikistan) and Maidanak (Uzbekistan) observatories (Meištas 1993; Meištas *et al.* 1999). The first WET run was successfully carried out at Mt. Maidanak in May 1992 and again in September 1992 (Meištas 1993; Meištas & Solheim 1993). At the end of the last millennium two Asian observatories – in Kavalur (India) and Beijing (China) – were ready to join the WET (Meištas *et al.* 1999) and Odessa National University Observatory (Ukraine) has arranged expeditions to Mt. Dushag-Eregdag Observatory (Turkmenistan) for possible participation of the WET campaigns (Dorokhova & Doroknov 2003). Recently, Lulin Observatory (Taiwan) has started to join the WET activities and participated in the latest two campaign observations, XCov 23 and XCov 24. However, up to now a considerable incompleteness in the longitudinal coverage still exists, in particular in middle Asia (AKA 'Asian gaps'), which leads to imperfection in light curves (see e.g., Solheim 2003) and therefore to difficulties in time series analysis.

2. About the site

2.1 Location

Mount Maidanak Observatory (Uzbekistan) is located in Central-Middle Asia atop one of the isolated peaks of Pamir and Alay mountain ranges. It has a longitude of $4^{\text{h}}27^{\text{m}}35^{\text{s}}\text{E}$, latitude of $38^{\circ}41'18''\text{N}$ and an elevation of 2540 m. The nearest town is Kitab (40 km northwards) while the nearest larger cities, Samarkand and Karshi, are 100 km and 150 km away, respectively. In Fig. 1, the location of the observatory is shown with the main world observatories and is zoomed in the separate window.

2.2 Seeing conditions

The site seeing and other atmospheric conditions in respect of observational astronomy (AKA ‘astroclimate’) have been investigated for more than thirty years. Site explorations indicated good astroclimatic parameters which were confirmed later on by modern evaluations using ESO’s Differential Image Motion Monitor (DIMM, Sarazin & Roddier 1990) started in 1996 and Generalized Seeing Monitor (GSM, Martin *et al.* 1994). As an example, the seeing distribution obtained by DIMM during a 3-year duration is shown in Fig. 2 (top panel) with a median seeing of 0.69 arcsec. The bottom panel of Fig. 2 shows the cumulative seeing (solid line) in comparison with those

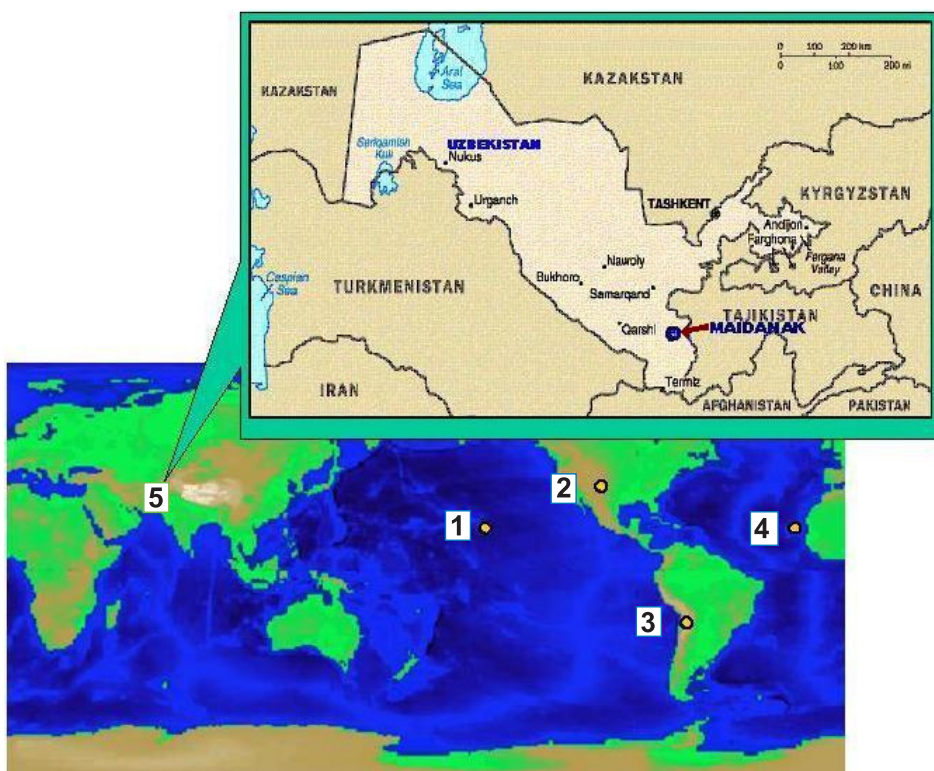


Figure 1. Global distribution of main good seeing optical observatories: (1) Hawaii, (2) NOAO, (3) Chile, (4) Canaries and (5) Maidanak.

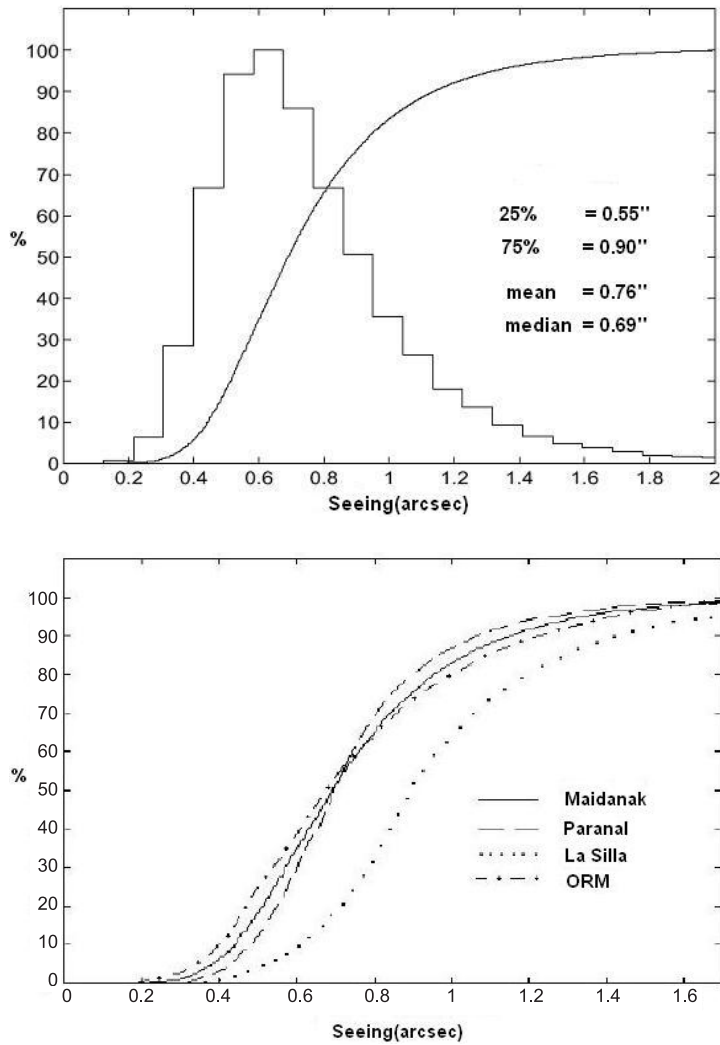


Figure 2. Mt. Maidanak seeing distribution (top panel) and its comparison with the sites in Chile and Canary islands.

Table 1. Some seeing parameters obtained by GSM (median values).

Outer scale (m)	25.9 ± 19.0
Seeing angle (arcsec)	0.67 ± 0.25
Isoplanatic patch (arcsec)	2.48 ± 0.70

of Chilean sites – La Silla (dotted line) and Paranal (dashed line) in 1989–1995 and Canarian site – Roque de los Muchachos Observatory (ORM) in 1996–1999 (Ehgamberdiev *et al.* 2000; Frogel 2002). The main seeing parameters obtained by GSM are given in Table 1.

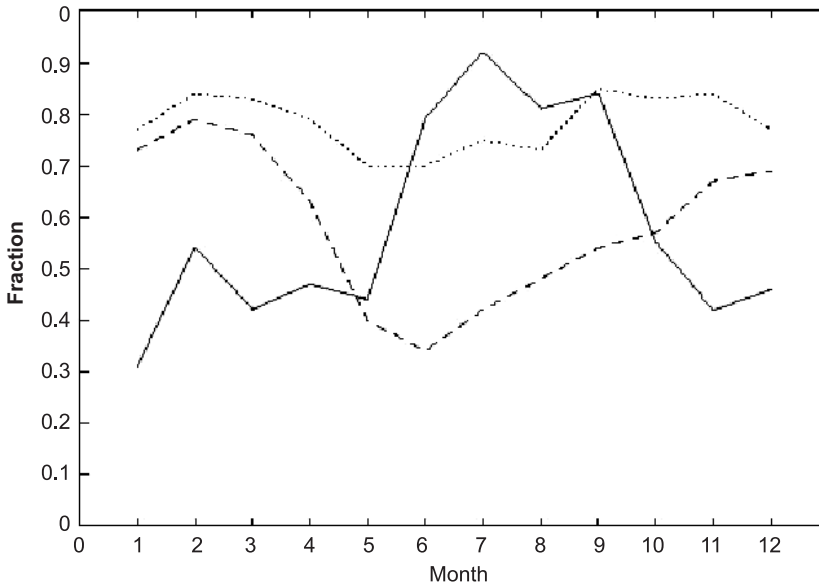


Figure 3. Clear nights distribution at Mt. Maidanak (solid line), La Silla (dashed line) and Paranal (dotted line).

In Fig. 3, we present monthly average fraction of clear nights at Maidanak in 1979–1985 (solid line) compared to the statistics of photometric nights at La Silla (dashed line) and Paranal (dotted line) in 1983–1997. Yearly fractions of clear nights are 58 per cent for Mt. Maidanak, 59 per cent for La Silla, and 78 per cent for Paranal, respectively. The total clear night sky is about 1900 hours per year. For more details on astroclimate atop Maidanak we refer readers to Sarazin (1999) and Ehgamberdiev *et al.* (2000).

2.3 Facilities

The observing facilities include 10 small-aperture telescopes. The main are 1.5 m and 48 cm telescopes made in St. Petersburg (Russia), one 1 m and three 60 cm Carl Zeiss telescopes (Germany), and one 50 cm telescope made in Kharkov (Ukraine). The Maidanak One-meter Telescope (MOT) had already been used in previous WET campaigns in combination with a special portable double-channel PMT photometer (Meištas 1993; Meištas & Solheim 1993). The 1.5 m telescope is equipped with an LN cooled back-illuminated CCD camera BroCam SITE005 with 2000×800 pixels. There are also a few old, hand-made single-channel PMT photometers. The currently operating DIMM instrument mounted on a standard Celestron-11 telescope (Sarazin & Roddier 1990) provides continuous and homogeneous seeing estimation coming in real time.

3. MOT upgrade and its possible use for WET

The MOT is a Cassegrain–Ritchey–Chretien–Coude system telescope, with a 1016 mm f/13 primary mirror. A recent photograph of the MOT is shown in Fig. 4. It was built in 1978 by Carl Zeiss in Jena (East Germany – former GDR) and has been operational

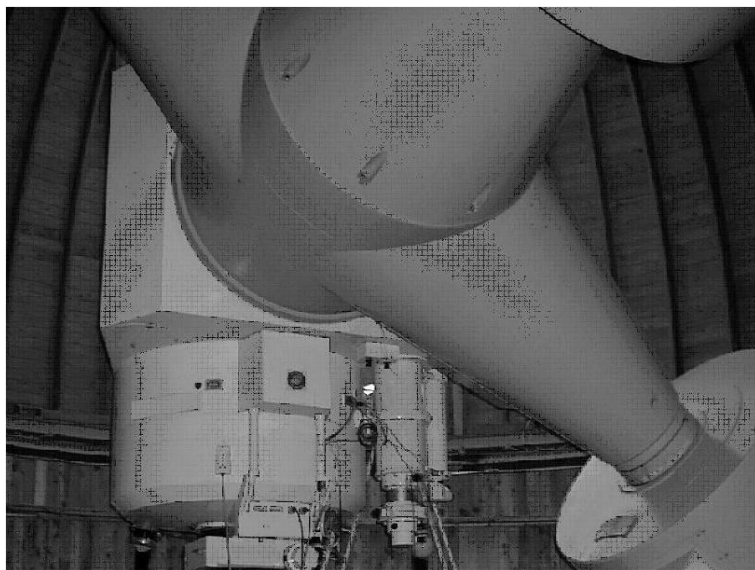


Figure 4. General view of the MOT.

since it first saw the light in 1981. In the summer and autumn (June–November) stellar photometry, mainly in the medium-band Vilnius photometric system, is conducted and in the winter some spectroscopy and radial velocity measurements are made (Meištas 1993). Naturally, the telescope electronics became old and outdated during this time thus making observing inefficient.

The idea to modernize and automate the MOT was initiated in 2001 when a team of devoted scientists from Institute of Astronomy of National Central University (Taiwan) observed at Maidanak (Hojaev *et al.* 2003). The financial resources were found later on when a group of astronomers from Taiwan, Lithuania, Latvia and Uzbekistan was funded by the Taiwan–Baltic Foundation to study open clusters. Simultaneously with conducting a joint research, the collaboration prepared and discussed possible renovation plans. The Yunnan Observatory group, which created a new computer control system (Zhou *et al.* 2001) and successfully upgraded their very similar telescope (the elder brother of the MOT – analogous 1 m RCC Zeiss telescope with the serial number differing from that of MOT by just 1!), agreed to join the collaboration and offered to fabricate the new computer control electronic components. A multilateral agreement has thus been signed. The installation of the new electronic system is planned for early 2005.

After the upgrade, the MOT should have computer control of pointing, tracking and focus, with display and analysis of telescope, instrument, and site status parameters. The observer can select a particular target from a graphical interface of allsky catalogs (e.g., GSC, DSS and USNO A or B) and the telescope will then slew (point) towards it. The new system also implements synchronization of the dome, *i.e.*, the dome slit follows where the telescope is pointing.

The CCD camera AP-10 (Apogee Instruments Inc., USA) is considered as the shakedown instrument for photometry and imaging of open clusters and star forming regions, and we plan to use it for the WET campaign. AP-10 has a front-illuminated Thompson Atmel THX7899 chip with specifications as presented in Table 2.

Table 2. Atmel THX7899 chip parameters.

Array size (pixels)	2048 × 2048
Pixel size	14 × 14 microns
Imaging area	28.7 mm × 28.7 mm
Pixel well capacity	270,000 e ⁻
Dynamic range	> 79 dB
CTE	0.99997
Peak QE	38 per cent (@ 720 nm)

Table 3. AP-10 camera system parameters.

PC interface	PCI controller card (proprietary)
Digital resolution	14 bits @ 1.3 MHz
Download time (typical)	5 seconds
System noise (typical)	13–25 e-RMS
Pixel binning	1 × 1 to 8 × 64 on-chip
Exposure time	30 milliseconds to 17 minutes (10 millisecond increments)
Frame sizes	Full frame, sub-frame, focus mode
Cooling	Thermoelectric cooler with forced air. Maximum cooling 35–38°C below ambient temperature
Dark current (nom.)	4 e ⁻ /pixel/sec (–10°C)
Temperature stability	±0.1°C
System gain	10–13 e ⁻ /ADU (14-bit)
Camera	Head Aluminum, hard black anodized. 7 inch × 7 inch × 2.5 inch
Weight	4 lb. 5.125 inch bolt circle for mounting. Optional Nikon, Canon or Thread mount
Operating environment temperature	–30° to 80°F
Relative humidity	10 to 90 per cent non-condensing
Power	Supplied by host PC. 25W max. power with shutter open and cooling maximum
Shutter	Melles Griot 63 mm iris
Remote triggering	TTL input to PC controller allows exposure start within 100 ns of trigger for applica- tions requiring precise synchronization of exposures to external events

The AP-10 camera system performance is presented in Table 3. As one can see from Tables 2 and 3, in spite of a relatively low QE of the camera especially in the blue range, it is quite large and at the same time fast (read-out time is 5 sec only). It should have possibilities for sub-framing and binning. The common shortcoming of

the Apogee camera is the problem with its bias. For AP-8E, for instance, the bias drift is smooth and slow within a night. In case of AP-10 we met with a sporadic jumping of the bias level. However the rest of the data are reliable and we simply excluded the 'jumped' (high-biased) frames from consideration.

One of the most important requirements of the WET observations is proper time-setting, *i.e.*, the correct fixing of the time of each measurement and synchronizing it with other observers and head-quarters. For this purpose we plan to use portable GPS station. It may be included as part of the new control system, as Yunnan Observatory has implemented for their 1 m (Zhou *et al.* 2001). Alternatively, voice time synchronization via a telephone will be used. The second item is operative link with WET head-quarters. Unfortunately, Maidanak has no internet link or reliable phone communication yet. But there were a few examples of successful use of the satellite communication systems. Bruvold (1993) describes the possibilities to use the micro-satellite communication with remote sites. During the WET runs in 1992 the TUBSAT-A (a student satellite of Technical University of Berlin launched as secondary payload of European Remote Sensing satellite ERS-1) was used which allowed sending and receiving short messages. However, it was not possible to transfer large volumes of data. Another successful usage of the satellite phone at Maidanak was by American amateur astronomers in 2003 during their visit to observe the Mars opposition when they had a voice communication with their partners in Tashkent and in the USA. Therefore the use of satellite communication systems for future WET campaigns is feasible.

One of the possibilities is a participation of Taiwan partners in the WET runs at Maidanak with such a device. Their immediate participation will be very desirable because they already have a good experience with WET activity at Lulin Observatory in Taiwan (XCov 23, XCov 24). We are also considering the use of a mobile phone for this purpose. The cellular communication systems as well as the internet are intensively developing also in Uzbekistan and becoming more spread. Obviously the signal will be weak in mountains far from settlements but hopefully we will be able to use the cellular phone with some signal amplifier to link with the internet and could use it at least for emailing. For instance, the cellular phones are in use in another mountain area north from Maidanak where the new large radio astronomy telescope should be installed soon.

4. Discussion and conclusions

The data presented in section 2.2 show that Maidanak has sky conditions rivaling those of the best sites worldwide. Its advantageous location can contribute to fill in the longitudinal gap in a global, continuous observing campaign. It can improve the WET coverage especially in cases of the monsoon season in India or be complementary to other network sites nearby. The recent distribution of observing sites which are involved in the WET activity is presented in Fig. 5. The WET community observatories as well as the Taiwan and Korea observatories that participated in the last campaigns are shown by filled circles while the Vainu Bappu Observatory in India is represented by the open circle. The Maidanak position is marked by a cross.

As mentioned earlier, the MOT has participated in previous WET observations (Solheim 2003; Meišt̃as *et al.* 1999 and references therein). A total of 4 successful WET campaigns have been carried out with the MOT (Meišt̃as *et al.* 1999). Figure 6

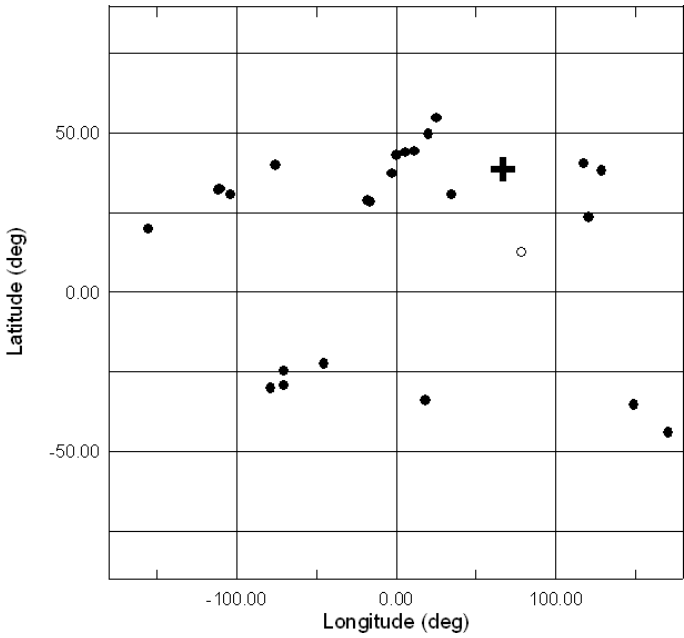


Figure 5. Distribution of the WET sites (filled circles); the open circle corresponds to the Vainu Bappu Observatory locus; the plus sign shows the location of the Maidanak Observatory.

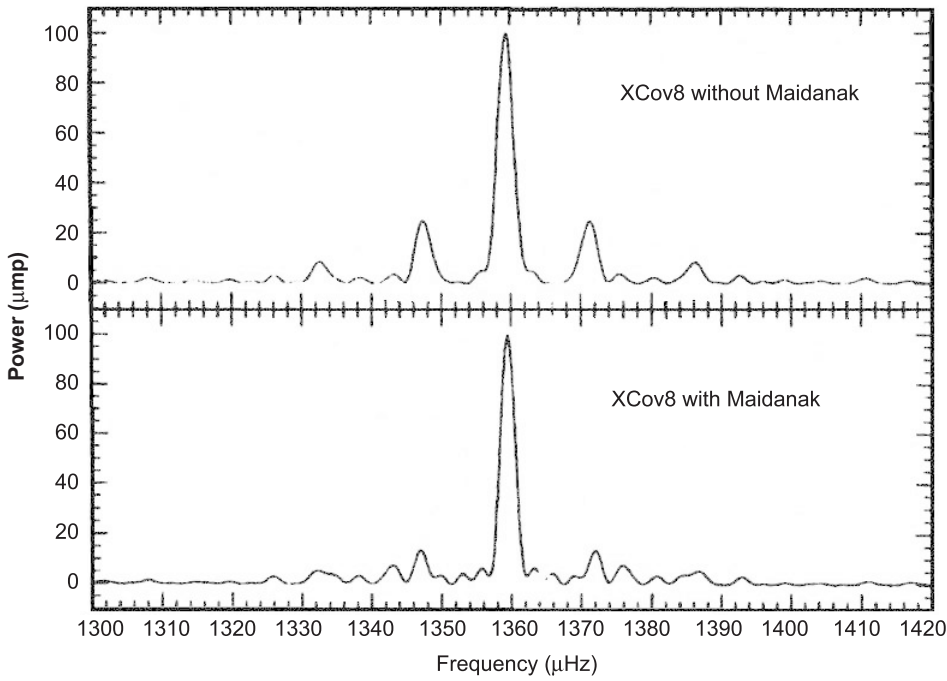


Figure 6. PG 2131+066: window function of Xcov 8 campaign in September 1992 with and without the Mt. Maidanak data.

shows the window functions without and with inclusion of the Maidanak data for PG 2131+066 reported by Solheim (1993). A more complete longitudinal coverage makes a cleaner power spectrum (Solheim 1993).

After the upgrade is completed the MOT should have a much higher observing efficiency and be more user friendly. With CCD cameras the photometry will be deeper and more precise than using, for example, the one-channel PMT photometer. Together with the favored geographical and sky conditions of the site, we expect Maidanak to become a useful part of the WET network.

Acknowledgements

The author thanks Prof. Wen-Ping Chen for his initiative to use MOT for the WET and for fruitful discussions. The author is also thankful to an anonymous referee for useful comments and suggestions. This work has been partly supported by grants from NSC Taiwan through the Taiwan–Baltic Foundation, and from STC Uzbekistan. Special gratitude goes to the Taiwan–Baltic Foundation for the upgrade project on the MOT. The author is also grateful for the financial support from the NSC for his visit to NCU.

References

- Bruvold, A. 1993, *Baltic Astronomy*, **2**, 492.
 Dorokhova, T. N., Dorokhov, N. I. 2003, *Baltic Astronomy*, **12**, 279.
 Ehgamberdiev, Sh. A. *et al.* 2000, *Astron. Astrophys. Suppl.*, **145**, 293.
 Frogel, J. A. 2002, *Memo for the SNAP project*, Lawrence Berkeley National Laboratory.
 Hojaev, A. S., Chen, W. P., Lee, H. T. 2003, *Astron. Astrophys. Transactions*, **22**, 799.
 Martin, F. *et al.* 1994, *Astron. Astrophys. Suppl.*, **108**, 173.
 Meiřtas, E. G. 1993, *Baltic Astronomy*, **2**, 479.
 Meiřtas, E. G. *et al.* 1999, *The 11-th European Workshop on White Dwarfs; ASP Conf. Ser.*, **169**, 176.
 Meiřtas, E. G., Solheim, J.-E. 1993, In *White Dwarfs: Advances in Observations and Theory*, Kluwer; NATO ASI, **403**, 549.
 Sarazin, M. 1999, URL <http://www.eso.org/gen-fac/pubs/astclim/>.
 Sarazin, M., Roddier, F. 1990, *Astron. Astrophys.*, **227**, 294.
 Solheim, J.-E. 1993, *Baltic Astronomy*, **2**, 382.
 Solheim, J.-E. 2003, In *The 6-th WET Workshop Proceedings*, 271.
 Zhou, J. G., Li, Z. L., Wang, B. 2001, *Publ. Yunnan. Obs.*, **2**, 42.

Argos: An Optimized Time-Series Photometer

Anjum S. Mukadam^{1*} & R. E. Nather^{**}

**Department of Astronomy, University of Washington, Seattle, WA 98195 1580, U.S.A.*

e-mail: anjum@astro.washington.edu

***Department of Astronomy, University of Texas at Austin, Austin, TX 78712 0259, U.S.A.*

e-mail: nather@astro.as.utexas.edu

Abstract. We designed a prime focus CCD photometer, Argos, optimized for high speed time-series measurements of blue variables (Nather & Mukadam 2004) for the 2.1 m telescope at McDonald Observatory. Lack of any intervening optics between the primary mirror and the CCD makes the instrument highly efficient. We measure an improvement in sensitivity by a factor of nine over the 3-channel PMT photometers used on the same telescope and for the same exposure time. The CCD frame transfer operation triggered by GPS synchronized pulses serves as an electronic shutter for the photometer. This minimizes the dead time between exposures, but more importantly, allows a precise control of the start and duration of the exposure. We expect the uncertainty in our timing to be less than 100 μ s.

Key words. Photometer—CCD—time-series—variable—instrument.

1. Requirements for an optimized time-series photometer

To study phenomena variable at short timescales, we use a technique called relative time-series photometry or high-speed differential photometry. A good time series photometer not only requires a precise measurement of the start time of an exposure, but also the duration of the exposure. Besides accuracy in timing, it must be able to provide sufficient time resolution to sample the variable phenomena well. These requirements must be additionally satisfied by a photometer of high quantum efficiency in order to qualify as an optimized time-series photometer.

For example, to study the hot ZZ Ceti stars that exhibit pulsation periods in the range 100–300 s, we need a suitable time resolution of 5–10 s. We would require a photometer that allows a short exposure time and introduces an insignificant dead time between consecutive exposures. Frame transfer CCDs are ideal for time-series photometry as they can provide contiguous exposures with no dead time. Photometers with a high readout time are incapable of good time-series photometry, even if they allow a 1 s exposure time. We show the light curve of a pulsating white dwarf acquired using the low resolution spectrograph in imaging mode on the 9.2 m Hobby–Eberly Telescope at McDonald Observatory (Fig. 1). We used 7 s exposures to match the instrument dead

¹Hubble Fellow.

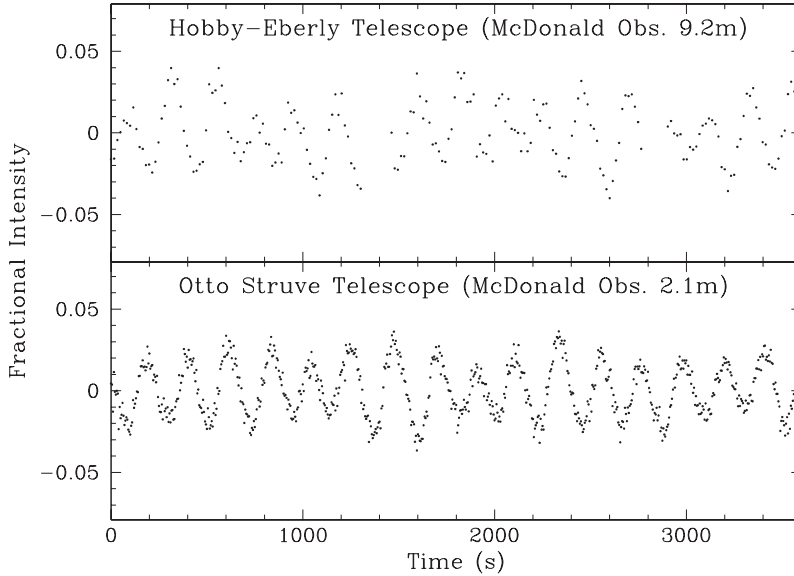


Figure 1. Light curves of a pulsating white dwarf acquired using the low resolution spectrograph on the 9.2 m Hobby–Eberly Telescope (top) exhibit uncertainties in period and phase larger by a factor of 1.5 than data obtained for the same time span using an optimized time series photometer on the 2.1 m telescope (bottom), also at McDonald Observatory.

time of 7 s. We also show observations of the same star using Argos (see section 2) with 5 s exposures on the 2.1 m Otto Struve telescope, also at McDonald Observatory. The data from the 9.2 m telescope exhibits uncertainties in period and phase larger by a factor of 1.5 than the 2.1 m data, acquired for the same time span (Mukadam 2004).

2. Argos: Designed for time-series photometry

We describe a high speed time-series CCD photometer, named Argos, designed for the prime focus of the 2.1 m telescope at McDonald Observatory (Nather & Mukadam 2004; see Fig. 2). One basic goal in the design of this instrument was to take advantage of the improved quantum efficiency that CCD detectors offer, so that we could obtain usable data on fainter stars than the Photo Multiplier Tube (PMT) instruments could measure, and data of better quality on those they could. We find that the CCD instrument is about nine times more sensitive than the PMT instruments used on the same telescope for the same exposure time (see Fig. 3). We can therefore find and measure variable white dwarf stars some 2.4 magnitudes fainter than the previous limit of magnitude 17.0, significantly increasing the number of such objects available for study.

A more long-term goal was to provide an instrument whose timing accuracy was high enough to allow a search for small deviations from the smooth secular change in pulsation frequency due to white dwarf cooling (e.g., Kepler *et al.* 2000), opening the possibility that such deviations, if found to be periodic, could demonstrate the presence of planet-sized objects in orbit around the white dwarf star (e.g., Kepler *et al.* 1991; Mukadam *et al.* 2001; Winget *et al.* 2003).

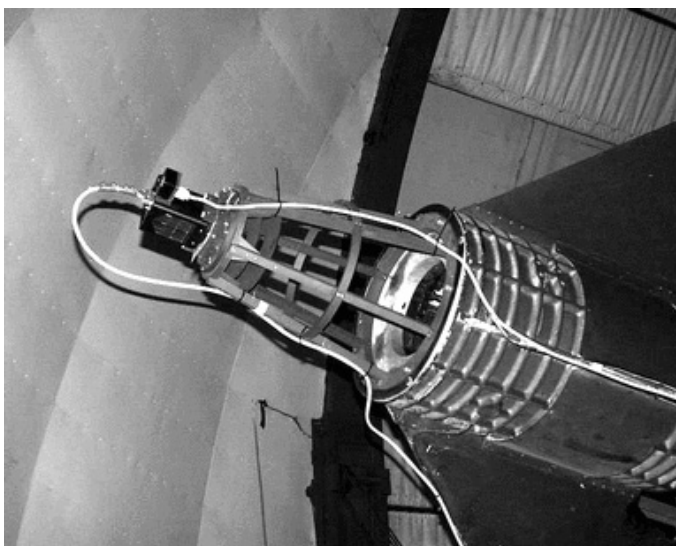


Figure 2. We show the CCD photometer Argos on its prime focus mount at the 2.1 m telescope at McDonald Observatory. Attached to Argos is a smaller CCD camera with a wide angle field of view (150°). It captures images of the dome slit at regular intervals, helping the user to decide when it is time to move the dome.

2.1 The CCD camera

Argos is based on a commercial instrument made by Roper Scientific, the Princeton Micromax 512 BFT NTE-CCD camera². We list the specifications in Table 1.

Table 1. Summary of camera specifications.

Active area	512×512 pixels, each pixel size $13\mu \times 13\mu$
Frame transfer time	310μ s
Readout time	0.28 s, unbinned full frame
Measured read noise	4 electrons RMS at 1 MHz (16 bit A/D conversion)
Gain	2 electrons/ADU
Dark noise	1–2 ADU/s/pixel at -45°C (thermoelectric cooling)
Quantum efficiency	30% at 3500 Å, 80% 4500–6500 Å, 40% at 9000 Å
Linearity	$\sim 1\%$ below 40,000 ADU (saturation at 65,000 ADU)

We acquire an image scale of 3.05 pixels per arcsecond (F/3.9) and a field of view of $2.8 \text{ arcmin} \times 2.8 \text{ arcmin}$ for our 512×512 pixel CCD chip. Frame transfer, initiated by pulses from a GPS system, allows us to obtain contiguous exposures as short as 1 s. The CCD is back-illuminated for improved blue sensitivity and provides a quantum efficiency of 80% in the wavelength range 4500–6500 Å. With thermoelectric cooling, we maintain the chip at a temperature of -45°C , and obtain a dark current of 1–2 ADU/s/pixel. The readout time for the entire chip with no binning is 0.28 s; the readout noise is less than 8 electrons RMS.

²<http://www.roperscientific.com/micromax.html>

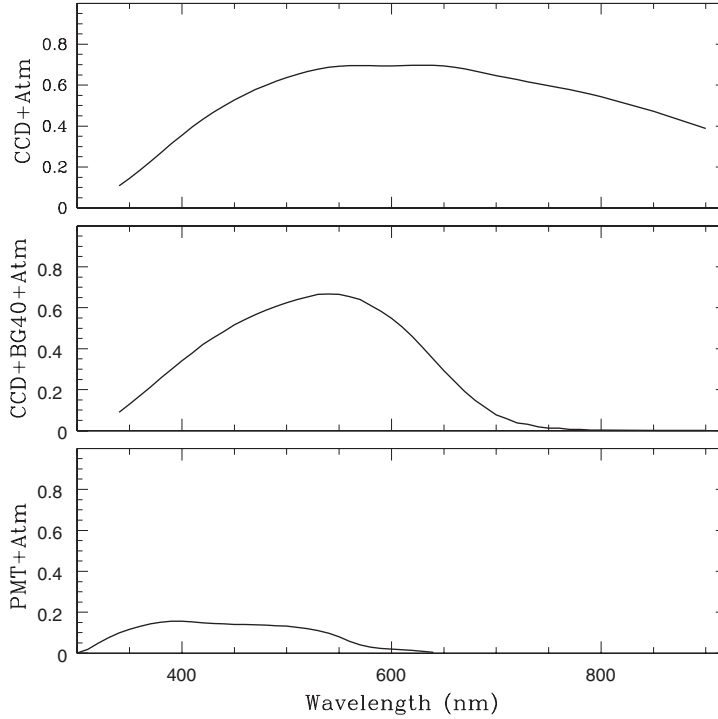


Figure 3. We show the wavelength response of the Argos CCD convolved with atmospheric extinction. The middle panel shows the modified response for a blue bandpass filter placed before the CCD. The lower panel shows the wavelength response of a PMT, a detector we used before Argos. The improvement in sensitivity of the Argos CCD vs. PMT is a factor of nine, as measured on the same telescope.

2.2 Wavelength response

We show the wavelength response of the CCD chip in the top panel of Fig. 3, taking atmospheric extinction into account (Mukadam 2004). We show the response of our instrument including a 1 mm BG40 Schott glass filter in the middle panel, while the bottom panel of Fig. 3 shows a PMT (R 647 Hamamatsu) wavelength response for comparison, convolved with atmospheric extinction. We used a PMT photometer, called P3Mudgee (Kleinman *et al.* 1996; Nather & Warner 1971), on the same telescope to gather data on pulsating white dwarfs before Argos.

2.3 Instrument timing

A time-series photometer must know precisely when an exposure started, and precisely how long it took. These are two different requirements, involving both a time epoch and an interval. The Argos timing system is based on a GPS clock designed primarily for precision timekeeping, although position information is also available. It consists of an oven-controlled crystal oscillator disciplined by signals obtained from a GPS receiver³. The time epoch is claimed to be uncertain by about 50 ns, considerably more precise than we need, but comforting.

³<http://www.trimble.com/thunderbolt.html>

We have assembled a simple count-down register on a small circuit card that can accept the 1 Hz timing pulses from the GPS clock, and can provide an output pulse to initiate frame transfer in the CCD camera (Nather & Mukadam 2004). The exposure intervals are thus contiguous, and determined directly from the clocking hardware. The jitter in the timing for this operation is hard to measure, but we expect it is $\leq 100 \mu\text{s}$. The timer card plugs into the parallel port on the camera control PC, so the countdown value (*i.e.*, the exposure time) can be set into it from software. Thereafter it operates independently, initiating frame transfer operations at the established timing intervals. Exposure times can be set to any integral number of seconds from 1 to 30.

Immediately following a 1 Hz timing pulse, the GPS clock provides information from which the precise epoch of the pulse can be determined. The information arrives encoded in packet form at the serial port of the PC at 9600 baud, where it can be read and unencoded by the software control program. The PC serial port is buffered so that the epoch (the time and date) for each pulse can be determined even if a second packet arrives before the first one is read by the program.

A second clock, somewhat less accurate, is available if the camera control PC is running the Network Time Protocol (NTP⁴) software, which obtains timing information over the internet by periodically contacting time servers and adjusting the PC system clock accordingly. It allows for internet time delays as well as it can, and averages the best readings it finds to keep the system clock in proper synchronization. The Argos control program relies primarily on the GPS clock for timing, but can use the NTP-disciplined system clock if the GPS time signals are not available. For observer assurance, it compares the time ticks from the two clocks, and displays their time difference in a status display window. After both clocks have been running for a few hours, the time difference is usually within a few milliseconds.

Note that although the camera originally arrived with a mechanical shutter, we removed it and chose to use the frame transfer operation as an electronic shutter. Mechanical shutters introduce a timing jitter, undesirable in a time-series photometer. Besides, they have a lifetime of an order of a million cycles and we do not expect them to last long with high-speed photometry.

2.4 The prime focus mount

We show Argos on its prime focus mount in Fig. 2, at the 2.1 m telescope. Argos mounts on the focus assembly, and the entire photometer moves along the optic axis when we focus the telescope. Design considerations for the prime focus mount necessitate that the optic axis passes through the center of the $6 \text{ mm} \times 6 \text{ mm}$ CCD chip, while the CCD plane remains perpendicular to the optic axis. We cannot rely on machining precision alone to center the optic axis on the chip, so we mount the camera on a plate that can be moved by half an inch in two orthogonal directions using x - y adjustment screws. These x - y axes align with the rows and columns of the CCD. We control the tip/tilt of the camera by a push-pull arrangement of screws, to orient the CCD perpendicular to the optic axis. We align the instrument every time the primary mirror is re-aluminized; Argos is sensitive to changes in the orientation of the optic axis larger than 24 arcseconds. When properly aligned, we expect the corners of the chip

⁴<http://www.ntp.org>

to show aberration from coma as a result of the parabolic primary mirror. This effect is calculated to expand a point image to about 1 arcsecond in diameter. We rarely experience sub-arcsecond seeing at McDonald Observatory, so we have not been able to verify this calculation.

2.5 Baffling Argos

Scattered light was initially a significant problem, and we chose hard (to corrode) black anodizing to eliminate the shiny aluminum surfaces of the mount. We also replaced the single crude baffle in the original design with a five-stage baffle system consisting of two thin baffles close to the camera, and three others in the body of the mount. The three baffles closest to the camera have square shaped apertures with rounded corners and are derived by projecting the light beam backwards from the CCD. Their apertures are a few per cent larger than the diameters of the light beam at those points. The other two mount baffles have circular apertures, which are 5–7% bigger than the light beam. The edges of all the light baffles are at an angle of 45° with respect to the optic axis to reflect light away from the CCD camera.

Our current images are flat to within a few per cent; the variations in the flat field come from structural non-uniformities in the CCD chip itself. This pattern is stable and can be removed, giving us residual variations of less than a per cent.

2.6 The camera control computer

The PC that controls the camera has fairly modest requirements by modern standards: it must have a PCI bus to accept the camera control card, a parallel port (for the timer card), a serial port (for the GPS packet information), enough memory to run the Linux operating system comfortably (256 MB is enough, but more is always better) and enough disk space to hold the images as they arrive (527 Kb each). Our current camera control PC runs a Pentium III at 1 GHz and is not pushed for time. The software prefers a display resolution of 1280×1024 so the various windows do not overlap each other. A 17-in LCD works fine. The PC also needs an ethernet card to connect to the internet, so the NTP software can discipline the system clock, and to receive pointing information from the computer that controls the 82-in telescope. This connection is also used to transfer image data to our Argos data archive in Austin (slowly), and to allow a remote login to run the camera for testing purposes (even more slowly).

We usually operate a second PC as well, with access to the disk on the camera control computer, so arriving image data can be examined by software not concerned with the data acquisition and recording process.

3. Controlling software

The control program is called Quilt 11 (q11), the most recent in a series of programs designed for time-series photometers (Nather & Mukadam 2004). The program also includes an online extraction routine, and displays the light curves in real time for the observer's perusal.

3.1 Program requirements

Once the camera has been set to operate in frame transfer mode, images arrive via direct memory access (DMA) at the end of each exposure and appear magically in memory. The program must first associate each image with its epoch (start time and date) before it is written to disk. This is not quite as straightforward as it sounds: a new exposure starts when a frame transfer operation finishes, so the epoch is available right away, but obviously the image is not—it's still being exposed. It only shows up after the next timing pulse (and its epoch) arrives, and then only after the readout process has finished. The program must account for this so that the proper epoch is associated with the appropriate image.

The data images are recorded in FITS⁵ format, with the epoch and other operating parameters in the header. Each image has its own file, so the file name must be generated automatically and the names must be sequential so they can be kept in proper time order. In our design the images arrive whenever a frame transfer pulse is generated by the clock—that is, all the time. We display the images using DS9.

Once data recording starts (in response to the “go” command), the extracted light curves are displayed by the DS9 plotting widget. The observer must then act as an aide to the program to do things it cannot do for itself: keep the star images on the chip, keep the dome out of the light path, and be prepared to shut things down if it rains. In addition to plotting the light curve, the program also prints columns of numbers on the terminal used to start the program: the image number, the elapsed time, and the extracted brightness for each of the marked stars with sky removed. The last column shows the average sky value that was subtracted from the target.

3.2 Software design

The q11 program is written in the C language, and consists of 5 separate executable processes in simultaneous execution. It is designed to run on a PC under the Linux operating system and to tolerate the presence of other programs running at the same time on the same CPU. Both incoming time and image data are buffered to maintain proper real-time operation. Details of its architecture and the on-line extraction algorithms are presented in Appendix A of our published paper (Nather & Mukadam 2004).

The program can also run in simulation mode, without a telescope, camera or GPS timing system. Previously recorded data images are read from disk into the same buffer used by the camera, and the timing is simulated from the system clock ticks. Users can use this mode to review data previously recorded.

3.3 Display

The data acquisition and on-line extraction process runs in a thread separate from the display process because of timing considerations: the display routines (DS9 and its plotting widget) are written in Tcl/Tk, an interpreted language, and are therefore slow. At the shortest exposure times the acquisition process can easily keep up but the display routines cannot. Arranged as separate threads of execution, they run in parallel, so acquisition gets its needed amount of CPU time even if display falls behind. Should

⁵http://fits.gsfc.nasa.gov/fits_intro.html

this happen, the user still sees the printed luminosity values appear right away, but the plotted values appear in clumps, rather than one at a time, whenever the plot widget gets updated. No data points are lost. Display of some of the incoming images may be skipped, but the DS9 window always shows the most recent one when it is updated. The q11 program can run under any window manager, but users notice how much more slowly the display windows are updated using Gnome or KDE, compared with WindowMaker, which is smaller and faster.

4. Characterizing Argos

The measured counts for a star contain statistical photon noise, scintillation noise mainly due to turbulence high in the atmosphere, seeing noise due to turbulence close to the telescope, and modulations from varying atmospheric transparency. Sky counts also include statistical fluctuations, modulations from changes in transparency, etc. Noise in the stellar and sky counts add in quadrature to give us the observed noise in the reduced light curve. We show the root mean square (RMS) scatter as a function of magnitude for light curves of constant stars, with exposure times of 10 s each, in Fig. 4 (Mukadam 2004). Such a plot can be used to derive the expected photometric precision for a star of given magnitude and for a given exposure time by simple scaling. We determine the instrumental magnitude by comparing the counts of the constant star to the known white dwarf in the field. This method is not reliable for stars much brighter than the white dwarf.

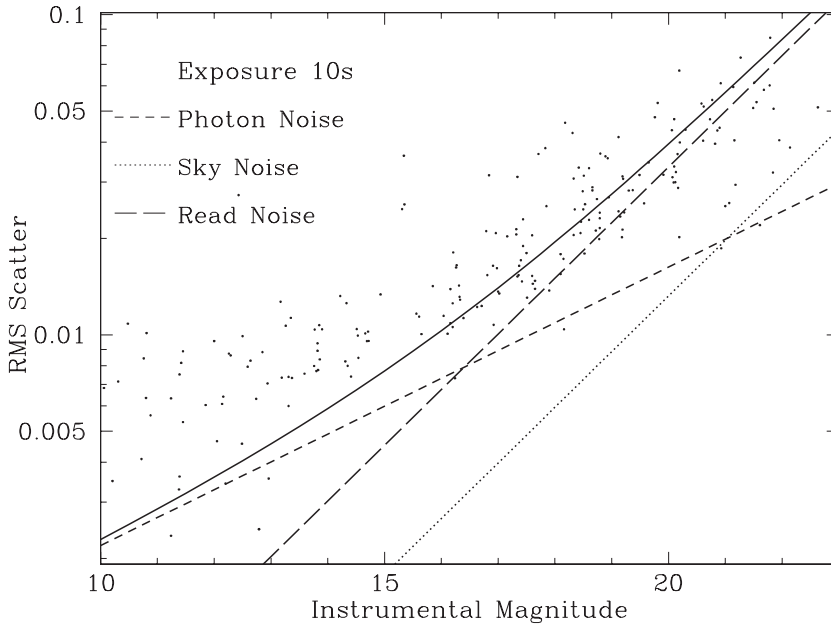


Figure 4. The RMS scatter in the light curves of constant stars shows photometric precision as a function of magnitude for 10 s exposures. Photon noise dominates the observed noise in bright stars. Noise in faint stars is governed by the quadrature sum of read noise and sky noise. For Argos on the 2.1 m telescope, we find this transition between “bright” and “faint” stars to be at 16.5 ± 1.0 .

Using the scintillation power spectrum published in Dravins *et al.* (1998), we estimate a 0.4% scatter in light curves of bright stars with exposure times around 1–3 s. We find the effects of low frequency scintillation and transparency changes mostly correlated (and hence removable) over the Argos CCD.

The scatter in Fig. 4 depends chiefly on the weather conditions, extinction, filters used, as well as magnitude and number of comparison stars in the field. Figure 4 shows a model of stochastic photon noise and sky noise, as well as read noise. Photon noise is the dominant source of noise for bright stars, while read noise and sky noise govern the floor for faint stars. The CCD specifications indicate a read noise of 8 electrons RMS, which would make the expected noise higher than the observed noise. By requiring that the quadrature sum of read noise and stochastic sky noise match the observed noise for faint stars, we measure a read noise of 4 electrons RMS. The magnitude at which we transition from a dominant photon noise to a combination of sky and read noise, helps characterize the sensitivity of the instrument. We determine this transition in the nature of noise to be at $g = 16.5 \pm 1.0$ for Argos on the prime focus of the 2.1 m telescope at McDonald Observatory (Mukadam 2004).

5. Results

Argos saw first light on 1st November 2001, at the 2.1 m telescope at McDonald Observatory, six months after we purchased the CCD camera from Roper Scientific. After struggling with initial problems concerning timing, scattered light, and so on, we slowly settled into an intensive program of observing 6 weeks every trimester. We discovered 35 new ZZ Ceti stars with Argos, mainly observing DAV candidates from the Sloan Digital Sky Survey (Mukadam *et al.* 2004). We discovered the faintest white dwarf pulsator known to date, a 20th magnitude DB variable, WD 0947+0155 (Nitta *et al.* 2005; in preparation). Argos was also instrumental in the discovery of optical bursts from a low mass X-ray binary MS 1603.6+2600 (Hynes *et al.* 2004), which confirmed that the compact object is a neutron star. We have achieved our long and short-term goals since the instrument has been in operation, and have established a list of target objects for the longer-term search for primordial planets.

The success story of Argos has also inspired a series of similar photometer designs at other sites such as Apache Point Observatory in the U.S.A., Indian Astronomical Observatory, and the Aryabhata Research Institute of Observational Sciences in India. An Argos-clone (called Bargas) was used successfully at Siding Spring Observatory, Australia, on the 1 m telescope in support of a Whole Earth Telescope (WET; Nather *et al.* 1990) run in May 2002.

Acknowledgements

We thank Dr. Frank Bash, Director of McDonald Observatory, for providing the funds to build Argos. We thank Denis Sullivan and Antonio Kanaan for their help in the commissioning of Argos. We also thank Antonio Kanaan for the use of his CCD data reduction routines.

References

- Dravins, D., Lindegren, L., Mezey, E., Young, A. T. 1998, *PASP*, **110**, 610.
- Hynes, R. I., Robinson, E. L., Jeffery, E. 2004, *ApJ*, **608**, L101.

- Kepler, S. O., Winget, D. E., Nather, R. E. *et al.* 1991, *ApJ*, **378**, L45.
Kepler, S. O., Mukadam, A., Winget, D. E. *et al.* 2000, *ApJ*, **534**, L185.
Kleinman, S. J., Nather, R. E., Phillips, T. 1996, *PASP*, **108**, 356.
Mukadam, A. S., Winget, D. E., Kepler, S. O. 2001, ASP Conf. Ser. 226: 12th European Workshop on White Dwarfs, 337.
Mukadam, A. S., Mullally, F., Nather, R. E. *et al.* 2004, *ApJ*, **607**, 982.
Mukadam, A. S. 2004, Ph.D. Thesis, University of Texas at Austin.
Nather, R. E., Warner, B. 1971, *MNRAS*, **152**, 209.
Nather, R. E., Winget, D. E., Clemens, J. C., Hansen, C. J., Hine, B. P. 1990, *ApJ*, **361**, 309.
Nather, R. E., Mukadam, A. S. 2004, *ApJ*, **605**, 846.
Winget, D. E., Cochran, W. D., Endl, M. *et al.* 2003, ASP Conf. Ser. 294: Scientific Frontiers in Research on Extrasolar Planets, **294**, 59.

High Resolution Stellar Spectroscopy with VBT Echelle Spectrometer

N. Kameswara Rao, S. Sriram, K. Jayakumar & F. Gabriel

Indian Institute of Astrophysics, Bangalore 560 034, India.

Abstract. The optical design and performance of the recently commissioned fiber fed echelle spectrometer of 2.34 meter Vainu Bappu Telescope are described. The use of it for stellar spectroscopic studies is discussed.

Key words. High resolution spectroscopy.

1. Introduction

An efficient high resolution spectrometer has been a long standing need at Vainu Bappu Observatory (that could cover a large region of the spectrum extending from 4000 Å to 1 μm in one exposure) for stellar spectroscopic programmes. The obvious choice that emerged is an echelle grating based spectrometer. The environmental conditions at VBO dictated the efficiency optimization to wavelengths longward of 4000 Å. The final choice was a spectrometer that has transmitting optics configured to operate in Littrow mode that provides a high resolution and a large wavelength coverage simultaneously by using an echelle grating at a large blaze angle (θ of 70°). The order overlap at the focus is separated by dispersing them in the perpendicular direction by the use of a cross dispersing prism. Prisms are preferred owing to their higher throughput efficiency and the fact that they give rise to a more uniform order spacing. The prism dispersion is greater in the blue where the free spectral range of the echelle grating (*i.e.*, the difference in wavelength between the centers of the successive orders) is less. The major advantage of this type of format is its suitability to image (or record) the spectrum on two-dimensional high quantum efficiency detectors, like CCD.

The conventional ‘coude’ scheme of bringing the star light to a laboratory based spectrometer for VBT requires seven reflections which may cause loss of about 10 per cent light at each reflection. It was realized that there would be much less light loss if transmitted by an optical fiber from the prime focus of the telescope directly to the spectrometer slit. The diameter of the fiber that satisfies the requirement of minimizing the sky light entering the slit while allowing most of the star light, is about 100 μm which corresponds to 2.7 arcsec at the prime focus. This size is slightly larger than the average seeing disc at VBT. There are other advantages as well for using a fiber to link the spectrometer to the telescope like minimizing the image guiding errors, image scrambling thus uniformly illuminating the slit (optics), etc. Thus the final choice was the use of a 45 metre long optical fiber to link the prime focus to the spectrograph slit. In the following account we describe the configuration of the echelle spectrometer. A more detailed account can be found in Rao *et al.* (2005).

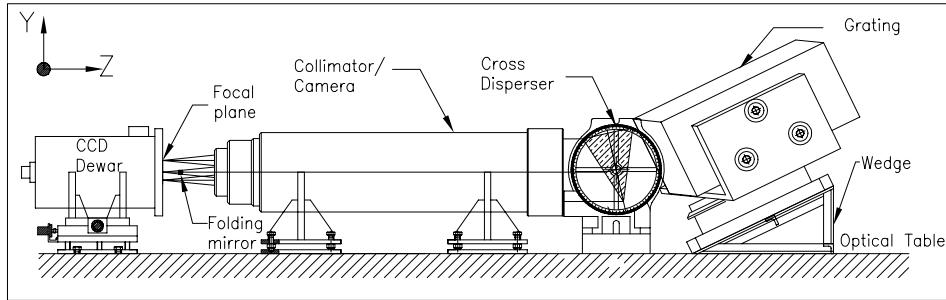


Figure 1. Opto-mechanical layout of the spectrometer. F/5 beam emerging from the slit is fed into the collimator by a folding mirror. The collimated beam then passes through the cross disperser (prism) and gets predispersed and illuminates the echelle grating. The dispersed beam from the grating again passes through the prism a second time and finally comes to focus on the CCD plane where the spectrum is recorded.

2. Design and optical layout

An echelle spectrometer like the conventional grating spectrometer essentially consists of a slit, collimator, grating, camera and detector. The additional element is a cross disperser prism or a grating. The fundamental consideration is the maximum spectral resolving power based on the astronomical motivation.

In our case we chose it to be about 70,000 (4 km s^{-1}). Once this is fixed other considerations about the echelle grating, the beam size, the camera, collimator focal length, the detector size, etc., follow. The questions that are considered here are:

- How to fit most of the spectrum in a given detector real estate (*i.e.*, the CCD size).
- The extent of the wavelength range (*i.e.*, number of orders).
- Overlaps (free spectral range, number of lines/mm) and order spacing such that the inter-order background is minimal (*i.e.*, cross-disperser).
- The intensity variations along the order as well as across the orders should be smooth and less steep (*i.e.*, broad blaze function) so that nearly the same signal to noise could be obtained across an order without saturating in one place and unable to register in other place.
- Efficiency should be optimal (*i.e.*, throughput should be high) and scattered light minimal.
- Good sharp lines without tilts and wings (in plane design).
- Stability and repeatability (environmental, mechanical stability, isolation from the telescope, temperature control and dome vibrations, etc.).
- Spectral format should be amendable for easy reductions and calibrations.
- Easy operations.

We finally preferred the all transmitting system for the spectrometer with the optical surfaces anti-reflection coated to optimize the throughput rather than reflecting cameras with central obscurations. Following other optimized throughput spectrometers like Sandiford at McDonald Observatory (McCarthy *et al.* 1993), we opted for Littrow configuration (*i.e.*, the collimator also acts as a camera; the input and output beams trace almost the same path). Again to have less steep light distribution along the orders we preferred to broaden the blaze function by using an echelle grating with more number of grooves/mm. The echelle grating selected has 52.67 gr/mm and blazed at 70° (R2.6).

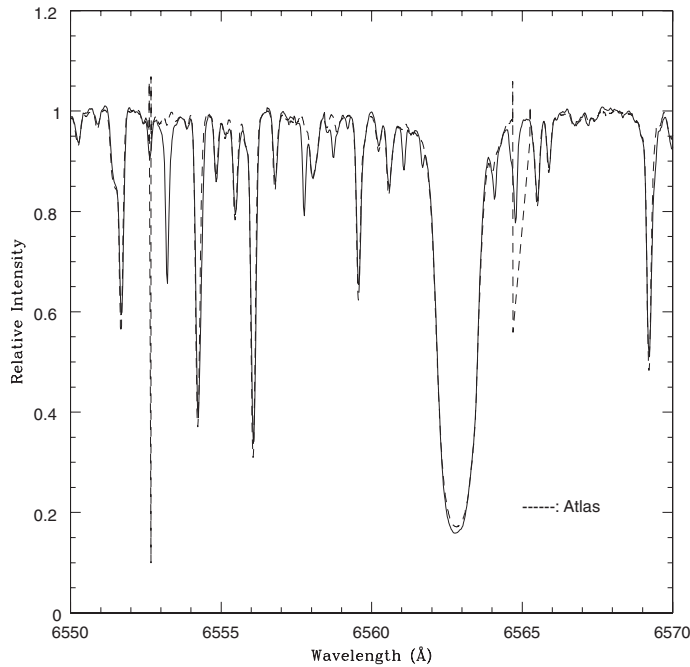


Figure 2. Spectrum of Arcturus obtained with fiber fed echelle spectrometer at VBT superposed on Kit Peak Atlas spectrum.

The expected resolution of the spectrometer with a slit width of $57\ \mu\text{m}$ and a collimated beam diameter of $151\ \text{mm}$ is 72,000.

The star light after hitting the primary mirror which comes to a focus as an $f/3.25$ beam, is fed to a $100\ \mu\text{m}$ core fiber located on the optical axis of the telescope. An image acquisition unit containing an intensified camera displays the stellar image and the fiber position on a monitor in the control room from where guiding is done. The light from calibration lamps – Thorium–Argon for wavelength calibration and Xenon and Tungsten lamps for flat field, is also fed to the fiber through the same unit. The fiber output is collimated and converted to an $f/5$ converging beam that focuses on to the slit. The image size at the slit is $154\ \mu\text{m}$. The fiber output can also be used without the slit with a lower resolution for maximum throughput. The light after the slit is reflected into the collimator by a $10\ \text{mm}$ folding mirror that sends the beam at an angle of 0.05° to the optical axis of the spectrometer.

The collimator is a six element $f/5$ system designed for a beam size of $151\ \text{mm}$ and has a focal length of $755\ \text{mm}$. The first element (also the last element in the double pass mode) corrects for off-axis feeding of the beam. The system further contains a doublet and a triplet, all anti-reflection coated and chromatically corrected to a wavelength range of $4000\ \text{\AA}$ to $1\ \mu\text{m}$.

All the elements in the collimator is rigidly mounted in a single cylindrical tube with a provision to flush them with dry nitrogen. The collimated beam then passes through an LF5 prism of about $165\ \text{mm}$ height and an apex angle of 40° which pre-disperses (in the cross dispersion direction) the light and sends it to the echelle grating. The echelle grating is illuminated by the incident beam fully. The grating is the single

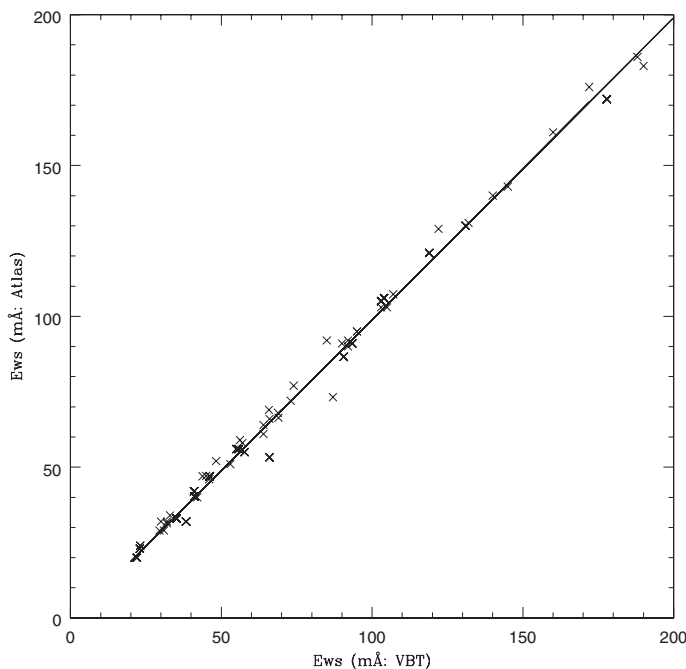


Figure 3. Comparison of equivalent widths in the spectra taken from VBT and Kitt Peak.

largest grating available and has a size of 408×208 mm. The dispersed beam retraces the same path and the prism cross-disperses for a second time (thus ensuring enough order separation) and enters the collimator for a second time. Now the collimator acts as a camera and focuses the dispersed spectrum on to a CCD chip. The emergent beam is again slightly off-axis and avoids the reflecting mirror that puts the beam in. The camera has a corrected field of about 60 mm diameter in the focal plane. The spectrum is recorded on a 2048×4096 pixels CCD camera exclusively built for the spectrometer. The CCD is a back illuminated, thinned Marconi chip with square pixels of $15 \mu\text{m}$ size and is placed in an LN2 cooled dewar. The controller and the associated electronics, and the dewar have all been built in the Indian Institute of Astrophysics laboratories. The slit size of $60 \mu\text{m}$ covers about four pixels on the CCD. The present CCD chip almost covers half of the area of the two dimensional spectrum format in the focal plane of the camera. The desired wavelength region is recorded either by moving the grating or CCD dewar. All the optical components are mounted on a Milles Griot vibration free table of size 8×4 feet. The opto-mechanical layout of the spectrometer is shown in Fig. 1. The CCD dewar is mounted on a stage with a provision to move in X–Y and Z directions, the direction of the optical axis of the collimator. The whole mechanical assembly is placed in an environmentally (temperature, humidity) controlled dark room – Coude laboratory.

It was realized early that a separate foundation might be required which is isolated from the rest of the telescope building and dome, for the coude laboratory. Thus four pillars have been incorporated in the building in the coude area exclusively for placing the spectrometer. The vibrational free table is placed on these independent foundations to isolate it from the building vibrations.

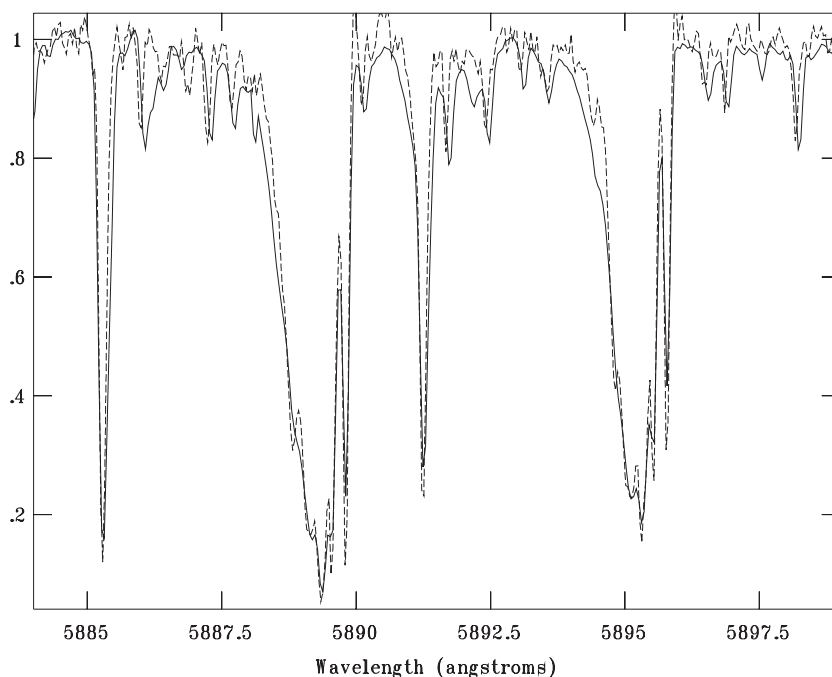


Figure 4. Comparison of the spectra of hydrogen deficient R CrB star, V854 Cen obtained with McDonald Observatory 107-inch coude echelle spectrometer with resolution 60,000 in February 1999 (full line) with the spectrum obtained by VBT fiber-fed echelle spectrometer on 13 Jan 2004 in Na ID line region.

All the operations of star acquisition, control of various lamp movements, etc., are done from the telescope control room remotely through a PC. The grating and prism movements could also be accomplished remotely through PC commands (encoder displays). The CCD stage movements also can be controlled remotely.

Presently the spectrometer operates in two modes with a 60 micron slit (resolution of 72,000) and without the slit (full fiber at 27,000 resolution) although there is a provision to change the slit width. By narrowing the slit we even obtained more than 100,000 resolution for HR 3117. However, there will be loss of light on the slit jaws whenever the slit is used till an image slicer is built for the system. Presently two other CCD systems are also available to be used with the spectrometer for programmes that do not require large wavelength coverage.

A 1024×1024 pixels of $24 \mu\text{m}$ size CCD system with a read out of 30s and a Pixcellent CCD system with 2048×2048 of $13.5 \mu\text{m}$ square pixels with the readout time of about 60s. Both the systems have higher quantum efficiency in the red and cover about the same wavelength range. For recording fast time varying spectroscopic phenomena these systems would be more suitable.

3. Some observational studies

In the last few months the spectrometer has been extensively used in the studies of stellar elemental abundances (Rao & Reddy 2005). The quality of the spectra obtained

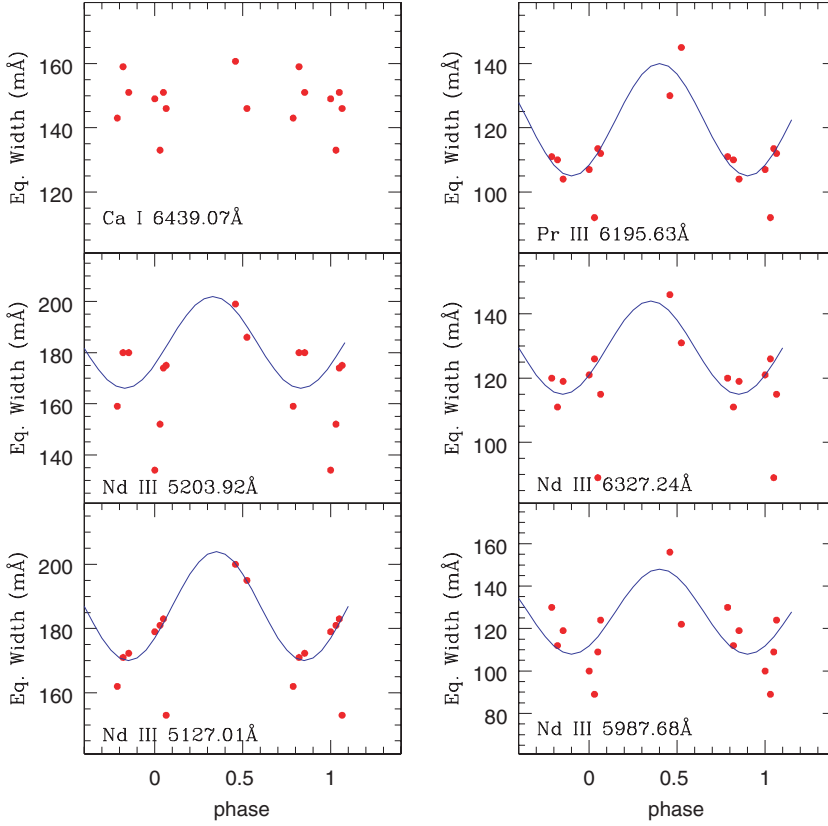
γ Eql (P = 12.21m, Cloudy 7 Oct 2004)

Figure 5. Variations of equivalent widths with period in γ eql. Variations are strong in rare earth lines where as Ca I lines do not show such variations.

can be illustrated from a comparison of Arcturus spectrum obtained with VBT and Kitt Peak Atlas (Hinkle *et al.* 2000). The Fig. 2 shows the spectrum of the star α -Boo (Arcturus) in the H α region obtained with VBT at 72,000 resolution (full line) compared with the spectrum displayed 'Visible and near-Infrared Atlas of the Arcturus spectrum' obtained at Kitt Peak at 120,000 resolution. Note the line to line matching and also the depth of the strong H α line, centered at 6563 Å, in both spectra match very well illustrating the high quality of the spectrum without any scattered light in the VBT spectrometer. Figure 3 shows the comparison of equivalent widths in the spectra to be an agreement within a standard deviation of less than 2 mÅ.

The spectrum illustrated in Fig. 4 is of the hydrogen deficient star V854 Cen that has been obtained with VBT echelle (dashed line) compared with the spectrum obtained with 107 inch telescope, McDonald Observatory, in 1999 at a resolution of 60,000 (full line) in the Na ID line region. The spectra are matched to the interstellar components. Note the narrowness of the interstellar lines and terrestrial water vapour lines in VBT spectrum compared to McDonald spectrum due to higher spectral resolution. However the central intensity of the stellar line blend matches in both spectra remarkably.

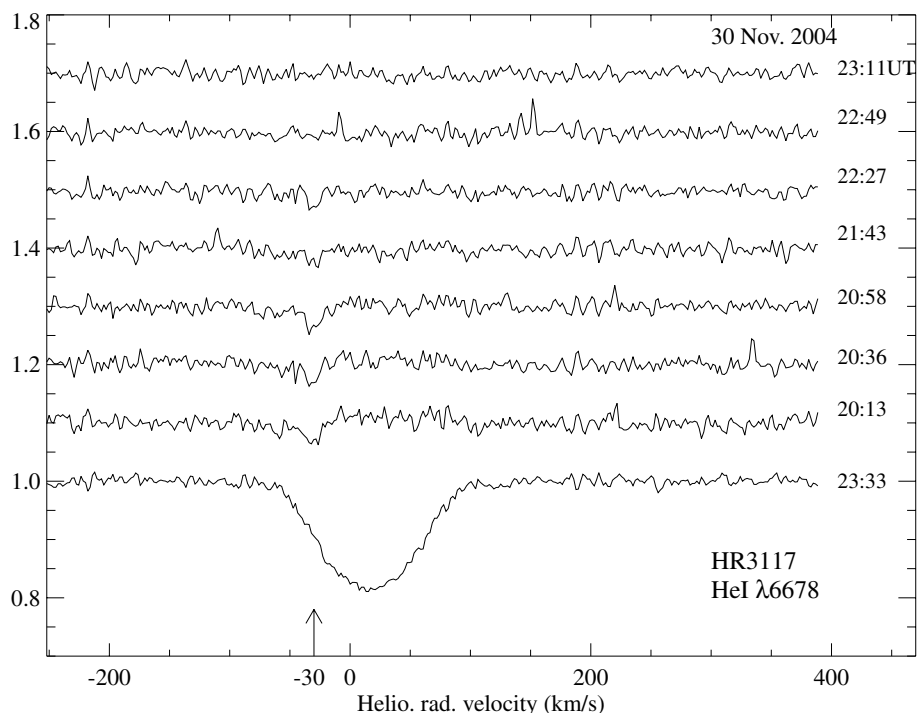


Figure 6. He I λ 6678 line and its variations with time in HR 3117. The divided profiles are arranged in increasing time sequence starting from the bottom. Note the extra absorption component at radial velocity -29 km s^{-1} that disappears with time.

In the context of the present workshop where variability is the central theme the following examples would illustrate the present capability of the instrument. The roAp star HR 1217 has been observed with a time resolution of 1 min at a spectral resolution of 28,000. The variation in the equivalent widths of a few selected lines is illustrated in the period of 12.2 min. Such variations are already well known in the star. The observation have been obtained using $1\text{K} \times 1\text{K}$ EEV CCD camera. More interesting are the observations of the suspected β Cephei star HR 3117 (χ Car). The star is supposed to show light variations with a period of 2 hours 25 minutes (Elst 1979). We have obtained spectra almost every 20 min on 30 November 2004 for a period of $3\frac{1}{2}$ hours continuously. The spectral resolution achieved is about 65,000 and the S/N ranged from about 108 to 180. During this period interesting changes have been seen in the He I 6678 Å line. As shown in the Fig. 6, the broad He I line extends from -68 km s^{-1} to 108 km s^{-1} at the base. The profile shows changes mainly on the blue side becoming from an asymmetrical line to a more symmetrical line. These changes are at 3 to 4 per cent level in relative intensity. The changes are better illustrated by dividing each profile with the last profile in the sequence (and relatively more symmetrical) and the resultant profiles are arranged in a time sequence along with the dividing profile in Fig. 6. It is clear that a narrow (relatively) absorption component that was present initially at radial velocity of -30 km s^{-1} (the stellar velocity is given as 19.4 km s^{-1}) slowly disappears in about 2 h 30 min. The equivalent width initially is 12 mÅ . Since the component already exists in our initial spectra and disappears in

$2\frac{1}{2}$ hours suggests the phenomenon, if periodic, would have a period much longer than the canonical light period. (Handler (private communication) says it is not considered as a β Cehei star any more because of lack of detectable light variations in Hipparchus measurements.) The star has been classified as peculiar by Hiltner *et al.* (1969), a B3 Si star, because of silicon line strengths. The helium lines have been studied by Leone & Lanzafame (1997) and Leone & Catanzaro (1998). Both these sets of authors obtain normal abundance of helium. Leone & Catanzaro (1998) determine abundances of other elements that are close to the values of main sequence B stars (including Si). Their spectra obtained on four nights in 1995 at a resolution of 13,000 showed no evidence of spectral variability. They even conclude that HR 3117 is not a peculiar star as reported in general catalogue of Ap and Am stars. Our present observations obtained at much higher resolution do show changes in He I lines. Additional observations are being obtained.

The above account describes the present capabilities of the spectrometer. Further developments in improving the throughput and more importantly improving the accuracy of radial velocity measurements are planned.

Acknowledgements

We would like to thank the organizers of this meeting for providing a nice platform for developing possible collaborations, particularly in the area of instrumentation. We would like to express our appreciation to several colleagues and friends, too numerous to list here, who either participated or offered valuable advice and helped us one way or other in the realizations of this major project. NKR would particularly like to thank Prof. Ramsagar for his kind hospitality during his visit to ARIES.

References

- Elst, E. W. 1979, *IAU Inf. Bull. Var. Stars*, 1562.
- Hiltner, W. A., Garrison, R. F., Schild, R. E. 1969, *Astrophys. J.*, **157**, 313.
- Hinkle, K., Wallace, L., Valenti, J., Harmer, D. 2000, *ASP Series on "Visible and Near-Infrared Atlas of the Arcturus Spectrum 3727–9300 Å"*, Kitt Peak National Observatory, Astron. Soc. of Pacific, San Francisco.
- Leone, F., Catanzaro, G. 1998, *Astron. Astrophys.*, **331**, 627.
- Leone, F., Lanzafame, A. L. 1997, *Astron. Astrophys.*, **320**, 893.
- McCarthy, J. Ki., Sandiford, B. A., Boyd, D., Booth, J. 1993, *Publ. Astron. Soc. Pac.*, **105**, 881.
- Rao, N. K., Reddy, B. E. 2005, *Mon. Not. R. Astron. Soc.*, **357**, 235.
- Rao, N. K., Sriram, S., Gabriel, F., Raghavendra Prasad, B., Samson, J. P. A., Jayakumar, K., Srinivasan, R., Mahesh, P. K., Giridhar, S. 2005, *Asian J. Phys.* (in press).

Facilities at ARIES for the Nainital–Cape Survey

Ram Sagar¹ & David L. Mary²

*Aryabhatta Research Institute of Observational Sciences (ARIES), Manora Peak,
Nainital 263 129, India.*

¹*e-mail: sagar@aries.ernet.in*

²*e-mail: dmary@aries.ernet.in*

Abstract. A collaborative programme searching for mmag pulsations in chemically peculiar stars in the northern hemisphere was initiated in 1997 between Nainital, India, and Cape Town, South Africa. It was therefore named as the **Nainital–Cape Survey** programme. The detection limits imposed by the observing conditions (including atmospheric noise and telescope size) at both Manora Peak and Devasthal sites are described. The scintillation noise on the best photometric nights is ≈ 0.1 to 0.2 mmag for these sites. Both places allow one to detect few mmag variation in bright stars ($B \leq 12$ mag), and are therefore particularly well-suited for carrying out the proposed survey work. The main characteristics of the three-channel photometer developed at ARIES for carrying out the observations are also presented. This excellent instrument has been used extensively since 1999 at the f/13 Cassegrain focus of ARIES' 104 cm telescope. In particular, it allowed the survey to result in the discovery of δ Scuti like pulsations in four Am stars, in one rapidly oscillating Ap star, and in a number of probable variables so far. The future prospects are then presented, which regard the acquisition of a high speed time series CCD photometer, a project to build a 3-metre class telescope at Devasthal, and collaborative observations with Indian and foreign astronomical sites.

Key words. Stars: chemically peculiar stars—oscillations— δ Scuti—asteroseismology.

1. Introduction

About 30 per cent of the stars in the classical instability strip of the HR diagram are δ Scuti stars which pulsate with long periods (tens of minutes or more) and amplitudes ranging from a few millimagnitude (mmag) to almost one magnitude. In contrast, some pulsating stars are rapidly oscillating Ap (roAp) stars. These roAp stars are cool, magnetic, chemically peculiar A-F IV-V stars which exhibit low-degree, high-overtone, non-radial p -mode pulsations with periods ranging from 5 to 21 minutes. The multi-periodic p -mode oscillations in roAp stars are of considerable significance because they allow the use of asteroseismology as a tool in the study of the chemically peculiar stars of the upper main sequence.

In 1997, out of the 31 roAp stars then known in the sky, only 3 were located in the northern hemisphere's sky. This large difference between the numbers of known roAp stars in the two hemispheres was attributed by Martinez *et al.* (2001) to the fact that greater survey efforts had been made at the South African Astronomical Observatory (SAAO), in comparison to other astronomical sites. This conjecture gave birth to the idea of a collaborative survey programme between both hemispheres. The aim of such a programme was to apply the proved and successful observational and data reduction techniques used in the Cape Survey by Martinez *et al.* (1991, 1994a, b) to a northern hemisphere observational site. The major requirements for such a site were that the sky conditions should be as good as at SAAO; the site should benefit from experienced astronomers in the area; and an important number of observing nights should be allocated to this programme on a one-meter class telescope. Since ARIES, Nainital fulfills all these criteria, it was decided in 1997 to initiate the so-called Nainital–Cape Survey programme. Further details regarding the survey programme and the procedures for observations and data reductions have been described by Joshi (2004) and references therein.

This paper analyzes the sites' characteristics, and exposes the current and future facilities dedicated by ARIES to this programme. The main characteristics and detection limits of the Nainital (Manora Peak) and the future Devasthal sites are analyzed in the next section. The three channel fast photometer, the future plans for the survey work, and the main conclusions derived from the results obtained so far are described in sections 3, 4 and 5 respectively.

2. Manora Peak and Devasthal as observational sites

The characteristic features of the atmospheric noise at both Manora Peak and Devasthal sites along with the detection limits imposed by them for a meter class optical telescope are given in the following sub-sections.

2.1 Sky transparency and scintillation

The presence of extremely low oscillation amplitudes (often < 1 mmag) in the survey work demand high photometric precision. There are two main sources of atmospheric noise that an observer has to contend with under photometric conditions. First, there are sky transparency variations that occur on a time-scale of 20 min or longer. The other source of atmospheric noise is scintillation. The latter sets, in practice, the lower limit of detectability of oscillation amplitudes in the survey stars – rather than photon statistics because all the candidate stars are bright, see section 2.2. Since the scintillation noise scales as $-2/3$ power of the telescope aperture, and as $3/2$ power of the air-mass (Young 1967), it is beneficial to use telescopes in the 1 m class or larger at relatively low air-mass.

Figure 1 shows a schematic amplitude spectrum of atmospheric noise consisting of both the sky transparency variations and scintillation, obtainable with 8 hours of observations on a photometric night at Manora Peak, Nainital with the 104 cm Sampurnanand telescope. It can be seen that the sky transparency variations have larger amplitude at $\nu \leq 1.0$ mHz. Its level drops down to the level of the scintillation noise somewhere between 1 and 1.5 mHz. On good nights, these sky transparency peaks are well separated from the frequencies of interest ($\nu \geq 1.0$ mHz) indicating that the roAp

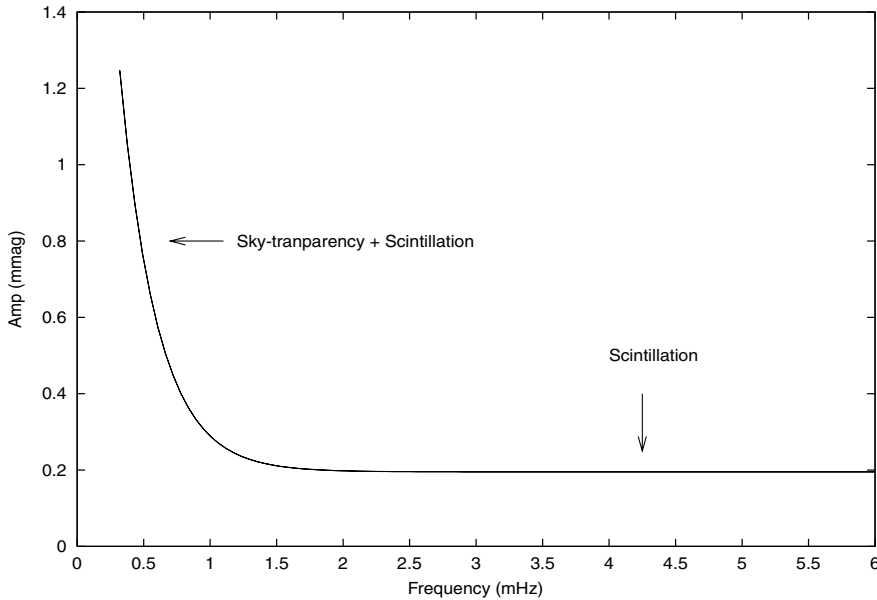


Figure 1. A schematic presentation of the sky-transparency and scintillation noise on a photometric night at Manora Peak, Nainital.

oscillations above the scintillation noise level can be detected without difficulty. On poor nights, the sky transparency noise can extend above 2.0 mHz making the detection of the roAp oscillations impossible. The level at which the sky transparency noise reaches the scintillation noise sets the lower frequency limit at which the oscillations can be detected and studied. Figure 1 indicates that the scintillation noise on a good photometric night is ~ 0.2 mmag. This is below the characteristic amplitude variations of most known roAp stars (a few mmag typically), and comparable to the scintillation level at the SAAO site.

Raw single-channel data of the variable star HD 98851 discovered at Nainital (Joshi 2004, 2005) are plotted in the upper panel of Fig. 2. One can notice gradual variations caused by changes in sky transparency, as well as oscillations due to the star. The application of a constant value of atmospheric extinction (0.26 mag/airmass) makes the star's oscillations clearly visible on a steady intensity level (Fig. 2, lower panel).

The results above show that Nainital is an excellent site to conduct the proposed survey work. Other characteristics of the site are reported by Kumar *et al.* (2000).

New optical observational facilities are currently being developed at another site, Devasthal, which is located about 50 km by road from Nainital towards the east. The results of an extended survey of the Devasthal site have been published by Sagar *et al.* (2000). Those indicate that the sky transparency at Devasthal is similar or better than that at Nainital, as it is located at a higher altitude, and situated further away from the light and other pollution sources. Consequently, Devasthal is a potentially excellent site for the present survey work as well.

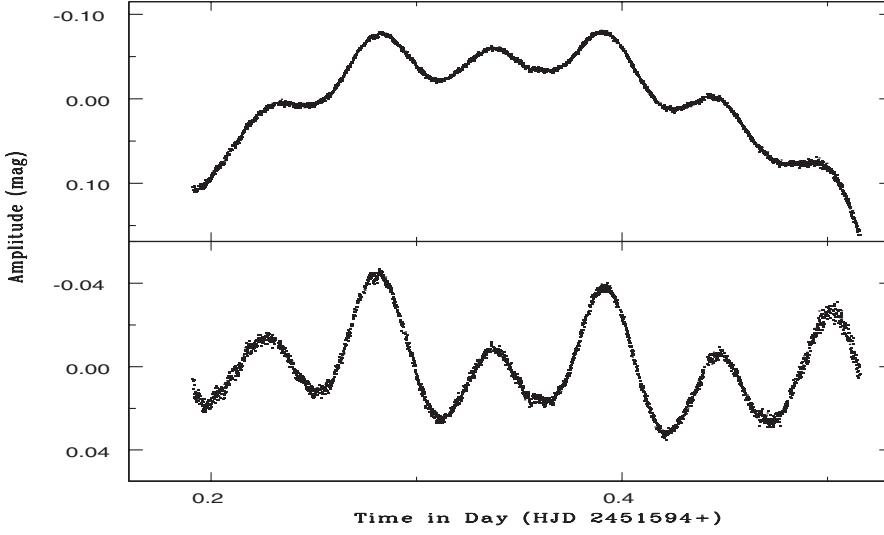


Figure 2. Light curve of HD 98851 obtained at Manora Peak. Upper panel: Raw data before extinction correction. Lower panel: Light curve after extinction correction.

2.2 Detection limits imposed by both atmosphere and optics

We present here an analysis of the detection limits determined by the scintillation noise, the photon noise and the background noise, to the detection of a variable signal. The study is made for both the Manora Peak and the Devasthal sites, and includes the effects of the optics (telescope + filters). The same telescope characteristics were assumed for both sites, though the one proposed to be placed at Devasthal will probably reveal superior optical performances to those assumed here.

Let us assume that a target star (magnitude m_λ , flux F_λ) pulsates with an amplitude of a few mmag (f). In differential photometry, the *signal* that we seek to detect is the *flux's variation* ΔF (corresponding to f) with respect to the average flux F_λ (corresponding to m_λ). This flux may be written as

$$\Delta F = F_\lambda(0) 10^{-0.4m_\lambda} (1 - 10^{-0.4f}), \quad (1)$$

where $F_\lambda(0)$ is the flux at magnitude zero (in $\text{erg/cm}^2/\text{s}/\text{\AA}$). Using an analysis similar to that of Mayya (1991), equation (6), the corresponding number of photons falling on the detector (just before amplification) during T seconds is

$$n_{\text{signal}} = \Delta F \Delta \lambda S \epsilon \eta 10^{(-0.4 x(\lambda) \sec Z)} T \left(\frac{hc}{\lambda} \right)^{-1}, \quad (2)$$

where $\Delta \lambda$, S , ϵ , η , $x(\lambda)$, $\sec Z$, h , and c are respectively the filter bandwidth (in $\text{\AA}$), the telescope's aperture (cm^2), the reflectivity of primary and secondary mirrors (dimensionless, 0.85 in our case), the efficiency of the system (including transmission of all filters, photometers' Fabry lens and quantum efficiency, dimensionless, 0.23 in our case), the extinction in the considered band (mag), the airmass, Planck's constant (erg s) and speed of light ($\text{\AA/s}$). The results below regard B band ($\lambda = 4380 \text{\AA}$) because

pulsations of roAp stars are generally monitored in this band. The values of all the atmospheric parameters for both sites can be found in Sagar *et al.* (2000) and Kumar *et al.* (2000).

The number of photons n_F corresponding to F , and the number of photons n_{sky} corresponding to the sky background can be derived in the same manner as in equation (2). (An aperture of 15 arcsec radius is used in the survey; this value was used to calculate the sky's background.)

The pulsation component oscillates around the average flux F . The average number of photons falling on the detector (just before amplification) is hence $n_F + n_{\text{sky}}$. Now, the joint effects of two phenomena will set the detection limit for a given site and telescope.

Firstly, the arrival of the photons is random, and in the absence of atmosphere, their statistics would be Poissonian, with mean and variance $n_F + n_{\text{sky}}$. However, because of random fluctuations of the air's refractive index occurring in the upper atmosphere, any deterministic flux propagating through these turbulent layers becomes random. The rms of this scintillation effect corresponds to a number of photons n_{sc} , which can be estimated from Dravins *et al.* (1998), equation (10). We obtain

$$\begin{aligned} n_{sc} &= 0.09 D^{-\frac{2}{3}} (\sec Z)^{1.75} e^{\frac{h}{h_0}} (2T)^{-\frac{1}{2}} (n_{\text{sky}} + n_F) \\ &= \alpha (n_{\text{sky}+n_F}). \end{aligned} \quad (3)$$

In the above equation, D is the telescope aperture (in cm), h the altitude (1951 m for Manora Peak, 2420 m for Devasthal) and $h_0 = 8000$ m. The integration time for the survey is $T = 10$ s.

Accounting now for the photon noise and the scintillation noise together, one sees that the photons falling on the amplifier follows a *doubly stochastic* process: the process is Poisson (because of random photon noise) with respect to a mean which is itself stochastic (because of scintillation). The study of these (so-called *mixed Poisson* processes) can be found mostly in the actuarial statistics literature (Grandel 1997); they have gained much attention recently in the field of astronomy, especially in extrasolar planet detection, see e.g., Canales & Cagigal (2001) and Aime & Soummer (2004). For such processes, the mean equals that of the Poisson process alone, and the variance is the sum of the variances of the two stochastic processes (photon noise and scintillation in our case). Consequently, using equation (3), the average number of photons falling on the amplifier is $n_{\text{sky}} + n_{n_F}$, and the associated variance is $n_{\text{sky}} + n_{n_F} + \alpha^2 (n_{\text{sky}} + n_{n_F})^2$. Hence, the overall (rms) noise, expressed in photons and including the joint effects of sky background, scintillation and photon noises becomes

$$n_{\text{noise}} = (n_{\text{sky}} + n_{n_F} + \alpha^2 (n_{\text{sky}} + n_{n_F})^2)^{\frac{1}{2}}. \quad (4)$$

This number can be evaluated for both Manora Peak and Devasthal sites. The signal-to-noise ratio (SNR) becomes

$$\frac{S}{N} = \frac{n_{\text{sig}}}{n_{\text{noise}}} = \frac{n_{\text{sig}}}{(n_{\text{sky}} + n_{n_F} + \alpha^2 (n_{\text{sky}} + n_{n_F})^2)^{\frac{1}{2}}}, \quad (5)$$

where the expressions of the involved noises are detailed above. The SNR tells us whether a variation of f mmag in a m_λ magnitude star is possible to detect or not.

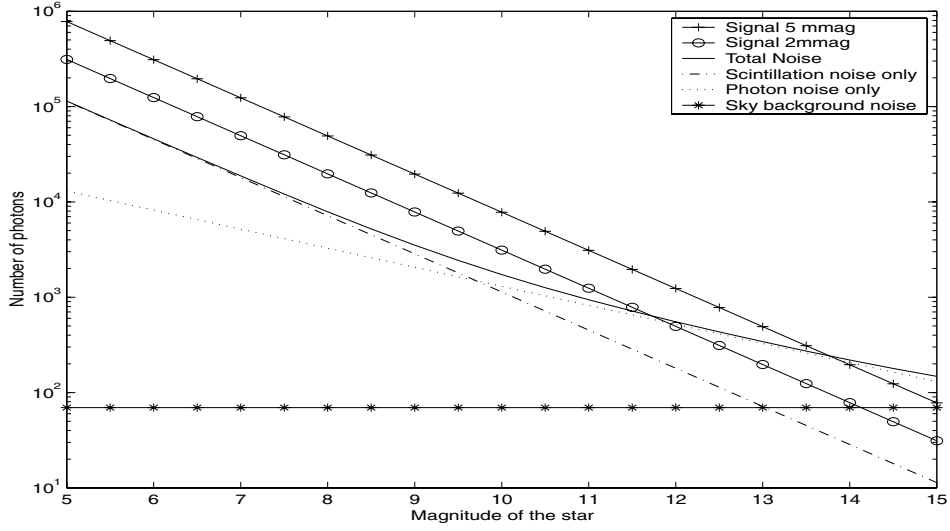


Figure 3. Signal and noise contributions at Devasthal for detecting a variation of 5 mmag and 2 mmag as a function of B mag for a star at the zenith.

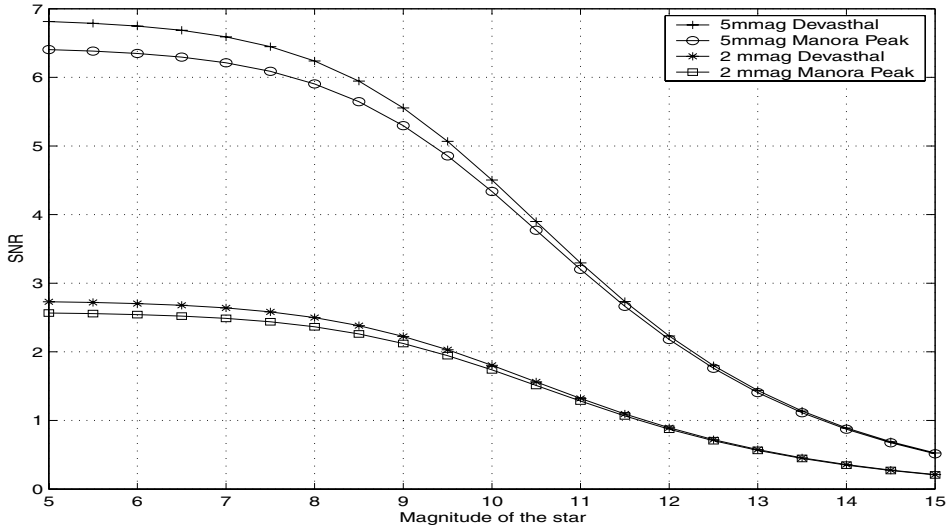


Figure 4. A comparison between the signal to noise ratios at Manora Peak and Devasthal for detecting 5 mmag and 2 mmag oscillations as a function of B mag for a star at the zenith.

Figure 3 illustrates the contributions of all the noise sources described above when trying to detect 5 mmag and 2 mmag variations as a function of B magnitude of a star at Devasthal. It can be seen that the sky background is negligible for all magnitudes of interest, and that the scintillation noise dominates the photon noise for bright stars.

Figure 4 compares the SNR (equation (5)) obtained at Devasthal and Manora Peak when trying to detect 5 and 2 mmag pulsations as a function of the B magnitude. Both sites allow the detection of 5 mmag pulsations for stars up to magnitude ≈ 14 in the

blue, and 2 mmag up to magnitude ≈ 12 . Also, one sees that Devasthal is even better than Manora Peak because of its more transparent sky and higher altitude.

The Devasthal site survey and the above results indicate remarkable photometric quality and stability of the sky conditions, which make the detection of a few mmag variations in bright stars ($B \leq 12.0$ mag) possible from Manora Peak, as well as from Devasthal for similar observing facilities.

3. ARIES' three-channel photometer

In order to carry out regular observations for the survey work at Manora Peak, Nainital, a new high speed three channel photometer was developed and fabricated (Ashoka *et al.* 2001; Gupta *et al.* 2001). Such photometers are essential for obtaining reliable and continuous time series data even under moderate sky conditions. The detector is a Hamamatsu R647-04 PMT made for photon counting applications at low light levels. The tube is blue sensitive with S-11 response. The photocathode has a size of 10 mm and is made of bialkali material. The tube has a gain of 2.2×10^6 . The quantum efficiency is 30% around 400 nm. Further details of the instrumental parameters are given by Ashoka *et al.* (2001).

The photometer is being used extensively for the survey work at Manora Peak since November 1999. We have mostly observed in B photometric passband as the tube response peaks in this passband. However, the photometer is capable of carrying out observations in U and V passbands also. While using three channels simultaneously, the sky contributions can be determined accurately even during variable moon-lit nights and dawn/dusk, because one channel of the photometer monitors the sky's contribution continuously. This also maximizes the utilization of the observing time on the telescope.

Regarding the sources of noises of the instrument, the high voltage (HV) supply has a capability of delivering 0.1 ma current, so that one unit itself can power three tubes. The low ripple (0.005% p-p) at maximum rating shows that the AC component is negligible, indicating that the HV is a perfect DC. The voltage is stable up to 0.01% for a change in 1V change in input, which is also negligible. This indicates that the voltage 1000v DC is very stable, and does not have any effect on the observed counts. Since the photocathode is blue sensitive, the tube does not need cooling. This is an end window tube with antimagnetic coating on the tube. This is the smallest (finger size) tube available for optical photometry. Because the tubes are small and uncooled, the whole photometer can be miniaturized. The dark counts at room temperature are ≈ 20 cps, which is again very low. Considering the overall performances on a 1 m size telescope in V band, one can observe a 16th magnitude star with a 10 s integration with instrumental SNR of 3.

Most of the time in a photometer however, the faintest star which we can observe depends on brightness of a star that one is able to see in the wide angle eyepiece, and on our ability to centre it in the diaphragm. Thus, the faintest stars may fail to be properly centred in the diaphragm, and one may thus lose a fraction of SNR – even though the corresponding flux could actually be detected by the instrument. In the case of CCD, we do not need to see and centre the star: we can centre the field and expose.

In conclusion, the global performances of the ARIES photometer are excellent for the purpose of this survey. Atmospheric scintillation, limited telescope aperture (and centering capabilities for fainter stars) are the major limiting factors for the discovery

and the variability measurements of pulsating chemically peculiar variable stars. This photometer has contributed to a major extent to the discovery of the new variables reported in Girish *et al.* (2001) and Joshi (2004, 2005).

4. Future facilities

There is a plan in ARIES to build a high speed time-series CCD photometer following successful measurements obtained with such instruments recently by Nather & Mukadam (2004), and Nather *et al.* (2005). As discussed above, CCDs will allow us to probe fainter objects in comparison to photomultiplier tubes; this in addition to a higher quantum efficiency will allow more precise photometric measurements.

Besides, the excellent review presented by Kurtz (2005) in this workshop has clearly shown that the pulsation effects in chemically peculiar stars are more prominent in some spectral lines than in the continuum, as they originate higher in the atmosphere, where oscillation effects are more pronounced. We plan to carry out such high-resolution spectroscopic observations using the facilities available in India (Rao 2005) as well as abroad (Mkrtychian 2005). We have also planned to build a 3 m class optical telescope, where high resolution spectroscopic observations will be routinely obtained for these chemically peculiar stars.

Finally, in order to have dense and long temporal coverage, we are currently organizing simultaneous observations from many observational sites in India and abroad, as a collaborative effort (Dorokhova & Dorokhov 2005; Hojaev & Chen 2005).

5. Discussions and conclusions

We have obtained precise photometric observations for over 150 chemically peculiar stars during 1997 to 2004 on about 100 photometric nights using the 104 cm Sampurnanand telescope of Manora Peak, Nainital. The remarkable characteristics of Manora Peak's astronomical site together with the use of successful and proved data reduction techniques developed at SAAO lead us to discover (only) one roAp star, but several δ Scuti, and other types of variables. Although the survey work is indeed on-going, the results obtained so far confirm an overwhelmingly large numbers of roAp stars in the southern hemisphere's sky in comparison to its northern counterpart. The discovery of only one roAp star in the presented survey so far *can not* be attributed anymore either to poor sky conditions at the site, or to lesser survey efforts. This difference in both hemispheres is thus extremely intriguing. As current observation do (and may be upcoming observations will) tend to confirm it, the need of a theoretical explanation for that phenomenon becomes more and more important. However, it should be noted that the availability of Michigan spectral catalogue (Houk & Cowley 1975 and later volumes) for chemically peculiar stars was a major advantage to the search for roAp stars in the southern hemisphere.

6. Acknowledgements

The financial assistance provided by the Department of Science and Technology, Govt. of India, New Delhi, for the development and fabrication of the three-channel fast photometer is acknowledged. The discussions held with collaborators from ARIES,

Nainital; ISRO, Bangalore; University of Central Lancashire, UK and SAAO, Cape Town, South Africa along with the numerous and essential information provided by Dr. Brijesh Kumar, Dr. Santosh Joshi and Dr. B. N. Ashoka to improve the quality of this paper are gratefully acknowledged. We are thankful to Dr. D. W. Kurtz for improving the paper as a referee.

References

- Aime, C., Soummer, R. 2004, *ApJ*, **612**, 85.
- Ashoka, B. N., Kumar Babu, V. C. *et al.* 2001, *J. Astrophys. Astr.*, **22**, 131.
- Canales, V. F., Cagigal, M. P. 2001, *Optics Letters*, **36**, 737.
- Dorokhova, T., Dorokhov, N. 2005, In: *Aries international workshop on Asteroseismology*, *J. Astrophys. Astr.*, this volume.
- Dravins, D., Lindegren, L., Mezey, E., Young, A. T. 1998, *PASP*, **110**, 610.
- Girish, V., Seetha, S., Martinez, P. *et al.* 2001, *A&A*, **380**, 142.
- Grandell, J. 1997, *Mixed Poisson Processes*, Monographs on Statistics and Applied Probability, Chapman & Hall.
- Gupta, S. K., Sagar, R., Joshi, S. *et al.* 2001, *BASI*, **29**, 479.
- Houk, N., Cowley, A. P. 1975, Michigan Spectral Catalogue of two-dimensional spectral types for the HD stars, Vol. 1, 2, 3 & 4, Dept. of Astronomy, University of Michigan.
- Hojaev, A. S., Chen, W. P. 2005, In: *Aries international workshop on Asteroseismology*, *J. Astrophys. Astr.*, this volume.
- Joshi, Santosh 2004, Ph.D. thesis, Kumaon University, Nainital.
- Joshi, Santosh 2005, In: *Aries international workshop on Asteroseismology*, *J. Astrophys. Astr.*, this volume.
- Kumar, B., Sagar, R., Rautela, B. S. *et al.* 2000, *BASI*, **28**, 675.
- Kurtz, D. W. 2005, In: *Aries international workshop on Asteroseismology*, *J. Astrophys. Astr.*, this volume.
- Martinez, P., Kurtz, D. W. 1994a, *MNRAS*, **271**, 118.
- Martinez, P., Kurtz, D. W. 1994b, *MNRAS*, **271**, 129.
- Martinez, P., Kurtz, D. W., Kauffmann, G. 1991, *MNRAS*, **250**, 666.
- Martinez, P., Kurtz, D. W., Ashoka, B. N. *et al.* 2001, *A&A*, **371**, 1048.
- Mkrtychian, D. E. 2005, In: *Aries international workshop on Asteroseismology*, *J. Astrophys. Astr.*, this volume.
- Mayya, Y. D. 1991, *J. Astrophys. Astr.*, **12**, 319.
- Nather, R. E., Mukadam, A. S. 2004, *ApJ*, **605**, 846.
- Nather, R. E., Mukadam, A. S., Winget, D. E., Sullivan, D. J. 2005, In: *Aries international workshop on Asteroseismology*, *J. Astrophys. Astr.*, this volume.
- Rao, N. K. 2005, In: *Aries international workshop on Asteroseismology*, *J. Astrophys. Astr.*, this volume.
- Sagar, R., Stalin, C. S., Pandey, A. K. *et al.* 2000, *A&AS*, **144**, 349.
- Young, A. T. 1967, *AJ*, **72**, 747.

The Concordia Station on the Antarctic Plateau: The Best Site on Earth for the 21st Century Astronomers

Eric Fossat

Laboratoire Universitaire d'Astrophysique de Nice, Université de Nice, France.

e-mail: Eric.Fossat@unice.fr

Abstract. On the Antarctic plateau, a joint project of French and Italian polar programmes is nearing completion: the Concordia station will be open for winter-over operation in 2005. The high altitude and high latitude of this site, the exceptionally cold, clear and stable atmosphere, its incredible astronomical seeing, the almost indefinitely flat snow surface and the not-so-difficult access make this site the most promising on Earth for future ground-based astronomical projects in various fields, including long term photometry, infrared high sensitivity imaging and high angular resolution and high contrast imaging.

Key words. Site testing—astronomical photometry—infrared astronomy—sub-mm astronomy—Antarctic astronomy.

1. Introduction

Twenty-five years ago, in November 1979, Gerard Grec, Eric Fossat and Martin Pomerantz started at the Amundsen–Scott South Pole station the first successful extended helioseismic campaign of history (Grec *et al.* 1980). This has since been regarded as the birth of helioseismology, a new science that required uninterrupted sequences of data sets significantly longer than 24 hr.

It was the possibility of non-stop, round-the-clock observations that attracted these pioneers. Later, several more campaigns were organized with similar goals, and an important infrared and sub-mm astronomical observatory has since developed in honour of Martin A. Pomerantz for his pioneering work in Antarctic astronomy. The local atmosphere at the pole is very cold, around -30°C during the less-than-3-month summer, and mostly between -50°C and -70°C in the winter, which lasts nearly 9 months. Moreover, this cold air is extremely dry – the driest desert climate on earth – with an amount of precipitable water pressure of the order of 0.25 mmHg in winter, and sometimes even less than 0.2 mmHg. These two parameters make the Antarctic skies very attractive for infrared and sub-mm astronomy, where new spectral windows can be accessed without the very costly space missions.

The very long polar nights and days, and the unique infrared possibilities are *obvious* reasons for developing Astronomy at the South Pole. There are other reasons, such as the clear sky statistics and the seeing. The dry climate on one hand and the absence of high altitude winds on the other (the extreme southern jet stream, named the polar vortex, rotates around the Antarctic continent and avoids the Pole itself) favoured



Figure 1. Starting in 2005 the two 3-storey towers of the Concordia station at Dome C will accommodate a small population of 16 winter-over scientists and technical staff. The 5-m high wooden Concordiastro platform, on the right side of this image, supports the site testing telescopes.

good expectations for these two parameters. Unfortunately, both have been somewhat disappointing. The Amundsen–Scott South Pole station is not far enough inside the continent to be protected from clouds, and, in fact, the clear sky statistics are not significantly better than 50%, the rest of the time the sky is covered in cirrus clouds. These clouds are at low altitude and are generally not totally opaque, but of course they reduce the possibility for very long measurements of any kind. The seeing itself has been measured in winter, with a disappointing mean result of the order of $1''.7$ (Marks *et al.* 1998). In fact, the balloon radio-soundings made in winter at the South Pole by these authors have explained this result. In Antarctica, the prevailing winds are catabatic winds, moving down from the high plateau to the coast and sometimes accelerated by the low pressure systems rotating around the continent (see Fig. 2). At the South Pole, these winds are not as fast as they can be on the coastal range, but they are present indeed, coming down from the Dome A area. They create a 200-m to 400-m thick layer of turbulence, present throughout the year that is responsible for the relatively poor seeing. At an altitude of 300 m above the surface, the seeing would be the best in the world.

2. The Concordia station at Dome C

At an altitude of 400 m higher than the South Pole itself (thus around 3280 m) and at a distance of about 1670 km from the pole (latitude 75° South, longitude 123° East), French and Italian Polar Institutes (IPEV and PNRA) are just finalizing the construction of the Concordia station (Fig. 1). It will accommodate a small population of 16 winter-over people starting in 2005. Originally, this site was selected by the glaciologists because it is one of the glacier domes where the horizontal motion of ice is nearly zero, a very favourable situation for drilling ice cores as ancient as possible.

As the prevailing winds are basically parallel to the ice motions (Fig. 2), the surface winds at Concordia should be significantly slower than at the South Pole. Indeed, as

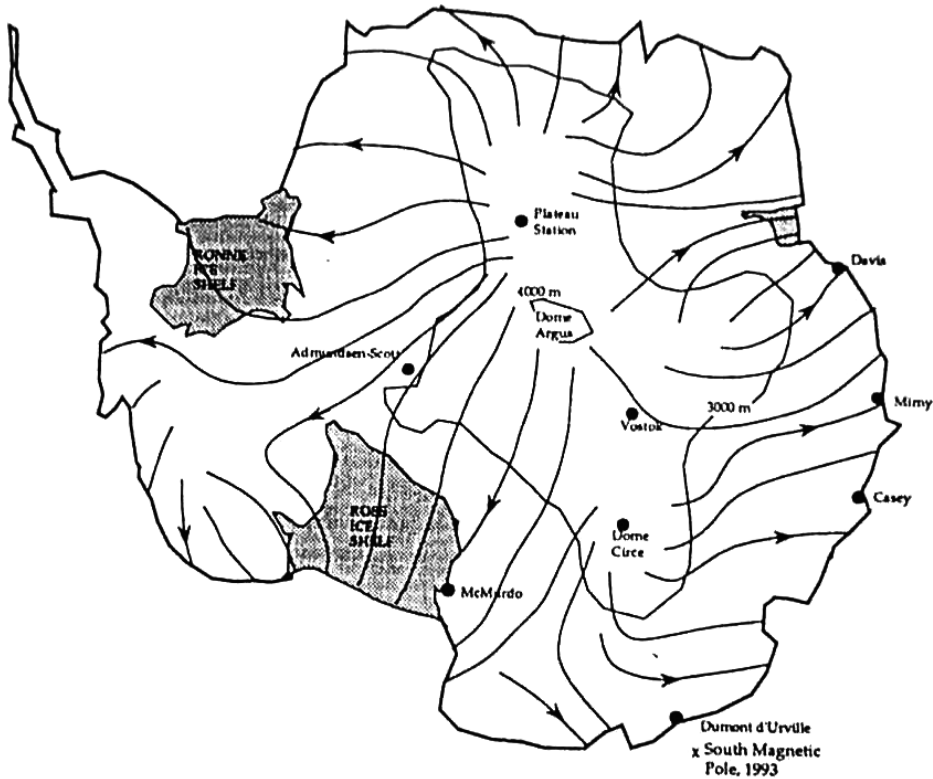


Figure 2. Prevailing winds in Antarctica, displaying the catabatic evidence (from Marks *et al.* 1998).

Fig. 2 clearly shows, Dome C is one of the high points of the polar plateau, and thus one of the starting points of catabatic winds. The result of this is an average surface wind of only 2.6 m s^{-1} , more than a factor 2 slower than at the South Pole.

The Concordia site is slightly drier and colder than the South Pole. It is then expected to be better for infrared and sub-mm astronomy. But being farther from the sea and less windy, it is also expected to be superior in clear sky statistics and, more importantly, in seeing quality, since the 200-m to 400-m turbulent layer of the South Pole is supposed to be absent at Dome C.

3. The Concordiastro and Aastino programmes

As soon as the station construction was started, LUAN (*Laboratoire Universitaire d'Astrophysique de Nice*) astronomers took advantage of their Antarctic experience (3 helioseismology summer campaigns at South Pole 20 years ago, two site testing campaigns at South Pole and Dome C about 10 years ago), and their site testing expertise. They proposed a 5-year site testing programme for Dome C, under the name of Concordia–Astro that was shortened into Concordiastro. It was proposed as a cooperative project with the *Observatoire de la Cote d'Azur* for the photometric part and *Observatorio Astronomico di Capodimonte* for the solar part. The aim is the determination of all atmospheric parameters that are relevant for astronomers. This

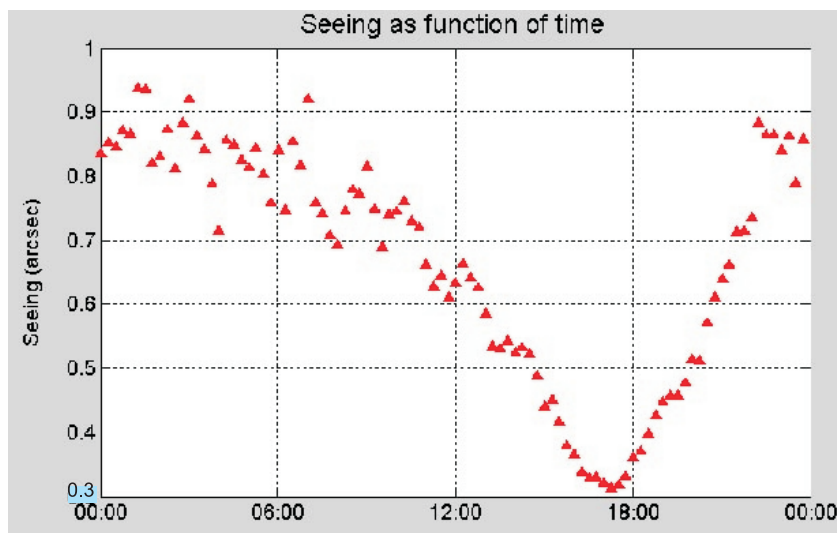


Figure 3. This is the 24-hr seeing measured by the DIMM instrument and averaged on the complete 2003/2004 summer season, about 2½ hours of very good quality each afternoon. Note that it offers to solar astronomers 6 hr every day of a seeing better than 0.5 and with a coronal sky most of the time! (from Aristidi *et al.* 2005).

programme has been developed in several iterations. At the time of the writing of this report, four summer campaigns have already been conducted, the fifth one is in progress, and the first winter-over should be starting in February 2005.

During the first summer campaign, a few meteorological balloons were launched, and the radio-soundings confirmed the absence of any fast wind at any altitude up to 25 km or sometimes 30 km. The second campaign was devoted to the launch of many more balloons, including a few tethered balloons for a more detailed analysis of the temperature gradients in the lower atmospheric layers (the first kilometre). A small 20-cm telescope was also tested for identifying the basic difficulties of operation in such a cold place. During the third campaign, the first wooden Concordiastro platform (Fig. 1) was erected and the first-ever seeing measurements with a DIMM (differential motion monitor) were successfully conducted on a bright star (Canopus) during the daytime on the bright light of the Antarctica glacier! The mean seeing was found, in daytime, to be much better than at the South Pole at night time; it is also better than in most solar observatories, but not outstanding though, a little over 1" (Aristidi *et al.* 2003).

But the next summer campaign used an improved telescope: the tube was made of invar and it was painted white. The mean seeing then was found to be very significantly better, less than 0.6" on average, with sequences around 0.1" (a number never seen anywhere else on Earth, even at night), and a systematic period of excellent seeing every day in the afternoon, local time – see Fig. 3 (Aristidi *et al.* 2005). During this fourth campaign, a second Concordiastro platform was also erected, the wooden “igloo” to be used at night as a local laboratory was constructed, and a second telescope was used for measuring the isoplanetic angle of the turbulence.

In parallel, the Australian group of University of New South Wales at Sydney, under the leadership of John Storey, installed in January 2003 the Aastino automated site testing observatory, and the two programmes have been organized in close cooperation

for an optimum efficiency and a minimum (but non-zero) redundancy. Aastino is a copy of Aasto, which has been operated at the South Pole since 1997–1998. They are both designed for automatic operation in wintertime, which needs to solve a power supply challenge that has not yet been fully solved all year round. However, Aastino was operated in 2003 from February to July and again in 2004 until May. It contains a SODAR (Sonic Radar), sensitive to the first kilometre of turbulence, and “summit”, a measurement of the sky transparency at $350\text{ }\mu\text{m}$. An automated webcam is linked to Sydney by satellite telephone, and a full sky camera, “icecam”, measures the cloudiness. During the 2004 season an additional instrument, Multi-Aperture Scintillation Sensor (MASS), was operated after February. It is a partial turbulence profiler (Kornilov *et al.* 2003) that measures scintillation through various apertures for reconstructing a rough vertical profile, with 5 independent values. MASS is sensitive to high atmospheric layers, while the SODAR is only sensitive to the lower layers, between 30 m and 900 m. The complementarity of these two sources of information has been tentatively exploited, with some inherent difficulty of calibration, for deducing winter seeing values of the whole atmosphere excluding the first 30 m (Lawrence *et al.* 2004). The mean value is estimated to be better than $0''.3$ during the nearly four months of data collected, a very promising number in good agreement with the best values obtained in afternoons during the summer daytime.

4. Some prospects

4.1 Site testing

The first winter-over campaign is about to start at the time of writing this report. Some night-time site testing results, then, still demand confirmation, as they are based upon extrapolations of the day-time measurements. There are, however, some indications that are already quite robust.

For instance, the clear sky statistics still need more investigation, but the preliminary results are extremely encouraging: the Aastino measurements give numbers of the order of 75% to 80% photometric, and our personal known record of a 5-d uninterrupted sequence of coronal sky during daytime at the South Pole was beaten by two sequences longer than two weeks each at Dome C (see Fig. 5 as an example of the midnight coronal sky, to be also regarded as very encouraging for solar physics).

The 300-m turbulent layer of the South Pole all year round is absent at Dome C (Aristidi *et al.* 2004). In summer, it is replaced by a changing temperature gradient layer of less than 100 m that strongly depends on the time of the day. The behaviour of the curve shown in Fig. 3 has been studied by means of many balloons launched on the same day, in order to check the variation of the temperature gradient. Figure 4 shows 10 samples of this gradient in the first 100 m. It happens that the best seeing occurs in the afternoon when the low-altitude gradient is essentially flat. On the other hand, the very best seeing values (around $0''.1$) were obtained very early in the season, when the ground temperature was still -50°C , or less. It now seems very likely that in winter only a thin layer of at most 30 m will represent a strong inversion layer, and should be stable most of the time, so that we have strong hope that the ground-based seeing will not be worse than the results estimated by the SODAR and MASS combination. Another indication is that through most of the winter time the SODAR itself does not see any significant signal, while it is sensitive to layers above 30 m.

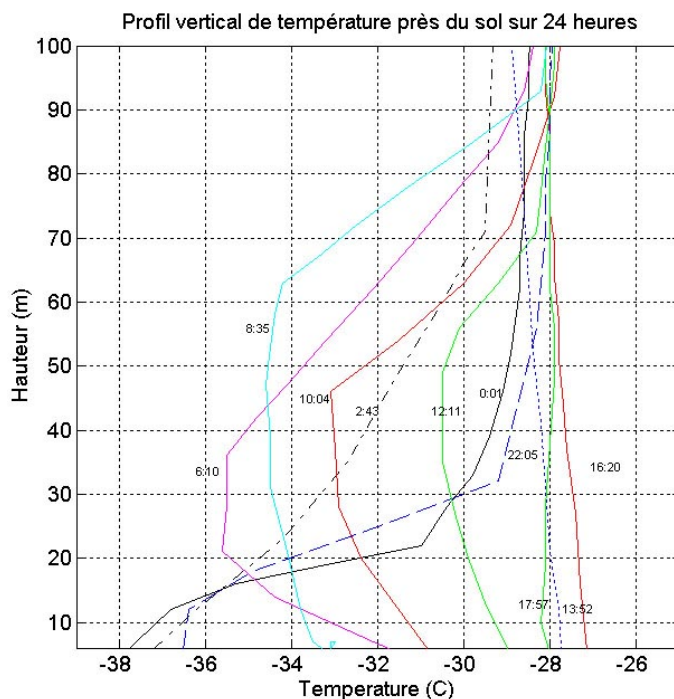


Figure 4. On a nice, clear day of January 2004, 12 balloons were launched in 24 hr for measuring the temperature gradient of the lower layers of the atmosphere. They show that when the seeing is best during the local afternoon; it corresponds to a very flat gradient on this plot. Therefore, the rest of the atmosphere, above a few tens of metres, is essentially never producing any significant turbulence since the seeing is always excellent during afternoons (from Aristidi *et al.* 2004).

The fourth summer campaign started very early in the season when the weather was closer to winter than to summer conditions. The high altitude wind was not absent during this first very cold week, with a first radio-sounding showing a 40 m s^{-1} wind speed at 20 km. The polar vortex, then, presumably does not always avoid the Dome C site in winter. However, as the best seeing values were obtained during that initial week, it shows that this fast, high-altitude wind speed does not significantly contribute to the seeing, although it may indeed result in an isoplanetic angle not as outstanding as expected. This was found around to be $8''$ on average. It is a factor 4 or 5 better than in other good sites. Winter measurements in July and August are now awaited.

Most of the (faint) winter-time residual turbulence is probably confined to the very low inversion layer. This short distance, combined with the slow wind speed, also results in an unusually slow turbulence. That will be one of the major targets of the night-time site seeing programme.

4.2 Infrared astronomy

The site will be better qualified by the end of 2005. But it is already quite clear that Dome C is the best site on Earth for its infrared background and transparency, for



Figure 5. This image, taken at local midnight on 22nd November 2003, with an ambient temperature of -51°C , shows the extreme “coronal” quality of the sky, as the Sun stands only 5° above the horizon (author’s image.)

its sub-mm new spectral windows and background stability, and for its seeing quality and time constants. The afternoon daily good seeing has already started to attract solar astronomers, and a first project of an infrared polarizing coronagraph has been proposed. At night, it seems that Dome C is the only place on Earth where deep infrared imaging, combining wide field and high angular resolution, can be envisioned at wavelengths longer than $2.4\ \mu\text{m}$. An extension of the already-achieved IR surveys to colder and darker objects (brown dwarfs, exo-planets), or also to more deeply hidden stars (young stars of low mass, or stars with thick dust envelopes for instance) can certainly be efficiently organized at this site.

4.3 Long-term uninterrupted photometry

Both sky transparency fluctuations and stellar scintillation are expected to be outstandingly less than anywhere else at night. With the high probability of extended periods of uninterrupted clear sky as long as two weeks and the geographical latitude that offers many stars permanently visible high in the sky, these parameters are obviously attractive for various stellar photometric programmes, such as the study of long-period variable stars, but also short-period and small-amplitude variables such as the solar-type oscillators. It is well known that the data interruptions constitute the major nightmare of all stellar photometrists. Another obvious very exciting prospect is the photometric search for exo-planet transits. Nobody has forgotten the transit of Venus across the

disc of the Sun in 2004. Similar transits in front of distant stars can be observed at Dome C, if thousands of stars are tracked permanently over weeks and even months.

4.4 *High angular resolution and high contrast imaging*

The faint and slow level of optical turbulence are strongly encouraging for the possibility of very high angular resolution imaging, especially at infrared wavelengths. Added to the exceptional photometric properties, this leads to the idea of direct detection of exo-planets, and possibly even of exo-earths. So far, all detections of planets orbiting around other stars have been indirect detections. Direct observations are indeed extremely demanding, as the planet is billions of times fainter than the star, and at an angular distance that is generally less than the broadening of the stellar image produced by the turbulence. In the infrared this factor of a billion can become only millions, and the turbulence is reduced. At Concordia this turbulence is already considerably reduced by the local atmosphere. Kilometric arrays of infrared telescopes being operated as imagers, very much like radio telescopes that have been exploited for decades already, can then be considered as realistic projects, and they are in the minds of those experts who think of the post-VLTI next generation of large-size interferometry. The scientific capabilities of such instruments are tremendous; they will go far beyond exo-planet detections, of course.

4.5 *The very early stages of the Universe*

Another exciting topic of modern astronomy deals with the polarization of the Cosmic Microwave Background (CMB), the fossil electromagnetic radiation coming from the very young Universe, around 300 000 yr after the Big Bang. Its anisotropy says something about the physics of this very beginning. This means Very High Energy Physics that cannot be reproduced in any laboratory. The polarization of the CMB is a trace left by the gravitational waves during the very early stages. Its amplitude is very tiny and its measurement requires extremely stable sky background in the microwave bands, something that the Concordia site is also very likely to be able to provide.

4.6 *Logistics*

Antarctica was, 56 yr ago, officially declared to be a continent totally devoted to science. This does not automatically mean that any scientific activity can easily be deployed there. Access to Dome C so far has only been possible during the 3-month summer season, and only by means of small airplanes operated by the French and Italian polar programmes. The Concordia station population is limited to 16 people, and not more than half of them can be scientists. The available electrical power is limited to about 200 kW, and a good fraction of it must be used for heating the buildings. Telecommunications are also very limited, so that a large part of scientific data exploitation will have to be done on the site, or wait for the next airplane that will transport data to more civilized places. On the other hand, and unlike most space observatories, the site is accessible every year. Heavy cargo loads can be transported by ground traverses from the Dumont d'Urville coast station, and these traverses have proved extremely efficient, reliable and also cost competitive. So, even if astronomical instruments must be made robotic for most of their exploitation,

they can also be permanently checked and repaired at least every year, and they can be technically updated every year, too, in order to use the most recent technological available developments. With the certainty of adaptive optics that will be much more efficient than anywhere else because of the atmospheric properties, many astronomers now dream of their next generation observatories at Dome C, since not everything will be done in space, of course. So far, the Concordia station has been a French–Italian programme. Its possible development as an important astronomical site will need a broadening of the international cooperation. We are just at the starting point.

Acknowledgements

The site testing campaigns of LUAN on the Concordia site have been totally supported, both science and logistics, by IPEV (Institut Paul Emile Victor). The author thanks D. Kurtz for translating in good English his first order approximation.

References

- Aristidi, E. *et al.* 2003, *A&A Letters*, **406**, L19.
Aristidi, E. *et al.* 2004, *A&A*, **430**, 739.
Aristidi, E. *et al.* 2005, in preparation.
Grec, G., Fossat, E., Pomerantz, M. A. 1980, *Nature*, **288**, 541.
Kornilov, V. *et al.* 2003, *Proc. SPIE*, **4839**, 837–845.
Lawrence, J. S., Ashley, M. C., Tokovinin, A., Travouillon, T. 2004, *Nature*, **431**, 278.
Marks, R. D., Vernin, J., Azouit, M., Manigault, J. F., Clevlin, C. 1998, *A&A Suppl. Ser.*, **134**, 161.

Unravelling the Nature of HD 81032 – A New RS CVn Binary

J. C. Pandey^{1,*}, K. P. Singh², S. A. Drake³ & R. Sagar^{1,**}

¹*Aryabhata Research Institute of Observational Sciences, Nainital 263 129, India.*

**e-mail: jeewan@aries.ernet.in*

***e-mail: sagar@aries.ernet.in*

²*Tata Institute of Fundamental Research, Mumbai 400 005, India.*

e-mail: singh@tifr.res.in

³*USRA & Code 662, NASA/Goddard Space Flight Center, Greenbelt, MD 20771, USA.*

e-mail: drake@olegacy.gsfc.nasa.gov

Received 2005 April 13; accepted 2005 November 28

Abstract. BVR photometric and quasi-simultaneous optical spectroscopic observations of the star HD 81032 have been carried out during the years 2000–2004. A photometric period of 18.802 ± 0.07 d has been detected for this star. A large group of spots with a migration period of 7.43 ± 0.07 years is inferred from the first three years of the data. H α and CaII H and K emissions from the star indicate high chromospheric activity. The available photometry in the BVRIJHK bands is consistent with the spectral type of K0 IV previously found for this star. We have also examined the spectral energy distribution of HD 81032 for the presence of an infrared colour excess using the 2MASS JHK and IRAS photometry, but found no significant excess in any band above the normal values expected for a star with this spectral type. We have also analyzed the X-ray emission properties of this star using data obtained by the ROSAT X-ray observatory during its All-Sky Survey phase. An X-ray flare of about 12 hours duration was detected during the two days of X-ray coverage obtained for this star. Its X-ray spectrum, while only containing 345 counts, is inconsistent with a single-temperature component solar-abundance coronal plasma model, but implies either the presence of two or more plasma components, non-solar abundances, or a combination of both of these properties. All of the above properties of HD 81032 suggest that it is a newly identified, evolved RS CVn binary.

Key words. Star–X-ray—star-variable—star-late-type—star-HD 81032.

1. Introduction

The presence of strong X-ray and nonthermal radio emission in late-type stars is a well-known indication of enhanced coronal activity (Drake *et al.* 1992; Güdel 2002). X-ray emission from coronae ($T_e \sim 10^6$ – 10^7 K) and chromospheric emission ($T_e \sim 10^4$ K) are also closely correlated, and thus stars with intense coronae will have strong chromospheric emission, as evidenced by their UV and CaII H and K line emission (e.g., Ayres

et al. 1995). While many active stars have been identified as such through their X-ray and radio emission, it is only through detailed optical photometric and spectroscopic studies that their activity can be classified into known types.

RS CVn binaries are perhaps the most common type of active star, at least for stars having spectral types in the range from late F through K. In an RS CVn binary, the two stars are usually tidally locked so that the rotational period of each star is approximately the same as the orbital period; however, the two stars are not undergoing mass transfer (i.e., they form a detached system). One star is generally a spectral type F to G dwarf or a subgiant that is ~ 1000 K hotter than its companion, which is usually a G to K type giant or sub-giant. In the more recent definition of the RS CVn class, there is no restriction on the spectral type of the secondary star or on the orbital period, except for the evolutionary constraint (Fekel *et al.* 1986). RS CVns have been further subdivided into three groups according to their orbital period: the short period ($P \leq 1$ d), classical ($1 \text{ d} \leq P \leq 14$ d), and long period ($P \geq 14$ d) groups. In the soft (0.1–2.0 keV) X-ray band, the luminosities of RS CVn binaries typically lie in the range of 10^{29} – 10^{32} erg s $^{-1}$ (Drake *et al.* 1989), compared to $10^{25.5}$ – $10^{29.5}$ erg s $^{-1}$ for normal late-type stars (Schmitt & Liefke 2004). VLA and ATCA observations at cm wavelengths have shown that the radio luminosities of RS CVn systems are in the range from $10^{14.5}$ to $10^{17.5}$ erg s $^{-1}$ Hz $^{-1}$ (Drake & Linsky 1986; Drake *et al.* 1989), which is also enhanced compared to that found for other late-type stars, typically $\leq 10^{15}$ erg s $^{-1}$ Hz $^{-1}$ (Güdel 2002).

The K0IV star HD 81032 (Houk & Smith-Moore 1988) was first noticed as an X-ray active star when it was identified as the likely optical counterpart of a soft X-ray (0.2–4.0 keV) source in the Einstein Slew Survey, 1ES 0920-13.6, with an X-ray flux of $2.1 \pm 0.9 \times 10^{-11}$ erg cm $^{-2}$ s $^{-1}$ (Elvis *et al.* 1992; Schachter *et al.* 1996). An X-ray source at this position was later detected in the ROSAT All-Sky-Survey (RASS) at a weaker (but statistically more significant) level of $3.9 \pm 0.4 \times 10^{-12}$ erg cm $^{-2}$ s $^{-1}$, named 1RXS J092253.7–134919 in the RASS Bright Source Catalogue (Voges *et al.* 1999). This star was detected on February 12th, 1993 as a radio source, with a 3.6 cm flux of 0.68 ± 0.05 mJy (Drake S. A., private communication). Given its spectral type and luminosity class, and its V-magnitude of 8.91 (Wright *et al.* 2003), the spectroscopic distance of HD 81032 is 140 ± 45 pc, implying an X-ray luminosity of $9.2 \pm 4.2 \times 10^{30}$ erg s $^{-1}$ (based on the RASS count rate and an assumed conversion factor of 6×10^{-12} erg cm $^{-2}$ PSPC cts $^{-1}$: a self-consistent X-ray luminosity based on the RASS data is calculated in section 8), and a radio luminosity of $1.6 \pm 0.7 \times 10^{16}$ erg s $^{-1}$ Hz $^{-1}$, where the distance uncertainty is the dominant contributor to the large errors in the luminosities. These values clearly show that the star HD 81032 has an active corona.

In this paper, we present extensive optical photometric and spectroscopic observations, as well as an analysis of archival X-ray data, of the star HD 81032. This is the first detailed optical photometric and spectroscopic study of this star, although a brief discussion of some of the optical photometry of this star was given in Pandey *et al.* (2002). Based on our long-term optical study, the star HD 81032 is shown to be a new member of the (long-period) RS CVn class.

The organization of this paper is as follows. The observations and data reduction are presented in section 2.. In section 3., we analyse the data for the periodicity. In section 4., we discuss the light curve and phase of minimum light. Chromospheric emission features are described in section 5.. In section 6. and section 7., the spectral

type and IR excess of HD 81032 are discussed, while the X-ray spectra are discussed in section 8.. Finally, section 9. summarizes the main results of this paper.

2. Observations and data reduction

2.1 Optical photometry

HD 81032 was observed in the Johnson *B*, *V* and Cousins *R* filter for 122 nights during four observing runs – year 2000–2001 (19 nights), year 2001–2002 (54 nights), year 2002–2003 (24 nights) and year 2003–2004 (25 nights) – at the Aryabhata Research Institute of Observational Sciences (ARIES). The observations were made with the 104-cm Sampurnanand telescope and using a $2k \times 2k$ CCD camera in years 2000–2003, and a $1k \times 1k$ CCD camera during the years 2003–2004. A few CCD frames were taken in the *B*, *V* and *R* filters on every night, with exposure times ranging from 2 to 60 secs, depending upon the seeing conditions and the filter used. Several bias and twilight flat frames were also taken during each observing run. Bias subtraction, flat fielding and aperture photometry were performed using IRAF¹. For the star HD 81032 the comparison and check stars were TYC 5471 1345 1 and USNO-A2.0 0750-06845737, respectively. Differential photometry, in the sense of variable minus the comparison star, was done, since program, comparison and check stars were all in the same CCD frame. UBVRI observations of HD 81032 along with some of the Landolt (1992) standard region SA 98 were obtained on 22nd February 2004 for photometric calibration. The average magnitudes of HD 81032 during the observing years 2003–2004 in the *U*, *B*, *V*, *R* and *I* filters were 10.242 ± 0.007 , 9.635 ± 0.005 , 8.614 ± 0.005 , 8.401 ± 0.003 , and 7.736 ± 0.003 , respectively.

2.2 Optical spectroscopy

Spectroscopic observations were carried out during 2003 January 20th to 24th at the Vainu Bappu Observatory, Kavalur with the OMR spectrograph fed by the 234-cm Vainu Bappu Telescope. The data were acquired with a 1024×1024 CCD camera of $24 \times 24 \mu\text{m}$ square pixel size covering a range of $1200 \text{ \AA}$ and having a dispersion of $1.25 \text{ \AA/pixel}$. The star was observed in the wavelength ranges of $3500\text{--}4700 \text{ \AA}$ and $5700\text{--}6900 \text{ \AA}$. Four spectra of HD 81032 in the $\text{H}\alpha$ region and three in the CaII H and K region were obtained. A signal-to-noise of ratio between 20 and 40 was achieved in these spectra. HD 71952, a K0 IV type star, was also observed as a reference star for HD 81032.

The spectra were extracted using the standard reduction procedures in the IRAF packages (bias subtraction, flat fielding, extraction of the spectrum and wavelength calibration using arc lamps). The spectral resolution was determined by using emission lines of arc lamps taken on the same nights. A spectral resolution ($\delta\lambda$) of $2.7 \text{ \AA}$ at $6300 \text{ \AA}$ and $3.7 \text{ \AA}$ at $4000 \text{ \AA}$ was achieved. All the spectra were normalized to the continuum and equivalent widths for the emission lines were computed using the IRAF task *splot*. The errors in the measurement of the equivalent widths of the emission lines were determined by measuring the equivalent widths of some moderate absorption

¹IRAF is distributed by National Optical Astronomy Observatories, USA.

lines for each spectra. We computed the standard deviation for the equivalent width of each feature, and finally determined the mean standard deviation to be 0.02 Å.

2.3 X-ray data

The star HD 81032 was observed and detected by the ROSAT PSPC detector during the ROSAT All-Sky-Survey (RASS) phase over a 2-day period from 1990 November 10th to 12th. The exposure time was 501s, and was accumulated in 26 separate short scans of this region of sky. The PSPC had an energy range from 0.1–2.4 keV with a (low) spectral resolution ($\Delta E/E \approx 0.42$ at 1 keV). A full description of the X-ray telescope and detectors can be found in Trümper (1983) and in Pfeffermann *et al.* (1987). The ROSAT X-ray data for HD 81032 were obtained from the public archives, the relevant RASS dataset being rs932025n00. Source spectra for HD 81032 were accumulated from on-source counts obtained from a circular region on the sky centered on the X-ray peak and having a radius of 3.85 arcmin. The background was accumulated from several neighbouring regions at nearly the same offset from the source.

3. Photometric light curves and period analysis

Photometric light curves corresponding to the four observing runs were taken. Figure 1 shows the *V* band differential light curves of the star HD 81032 at different epochs. We did not find any significant variations in the comparison star (see below). The yearly mean of the standard deviation (σ) between the different measures of comparison and check stars in the *B*, *V* and *R* filters was found to be 0.011, 0.01 and 0.01, respectively. Each light curve shown in Fig. 1 was analysed for periodicity. To find a period from unequally spaced data, we used the CLEAN algorithm (Roberts *et al.* 1987) in Starlink's PERIOD software. The power spectrum obtained using this method is shown in the inset of each panel of Fig. 1 along with the period determined. The CLEANed power spectra presented were obtained after 100 iterations with a loop gain of 0.1. The period is found to be constant within error for each epoch. To improve the period determination of the star HD 81032, the entire data from 2000–2004 were analysed using the same algorithm. Figure 2 shows the CLEANed power spectrum from the entire dataset. The highest peak in the CLEANed power spectrum corresponds to a period of 18.802 ± 0.074 d. The 18.802 d period is much more plausible than the 57 d period reported earlier (Pandey *et al.* 2002). The previous determination was mainly due to observational limitations as our early data were too sparse and highly uneven. Besides a subgiant like this one is unlikely to be synchronized in a binary of 57 days period.

4. Photometric variation and phase of minima

The Julian days of the observations were converted to the phases using the ephemeris:

$$\text{Phase}(\theta) = \text{JD } 2452307.761 + 18^{\text{d}}.802E$$

where the initial epoch corresponds to the conjunction with the first minimum observed. Figure 3 shows the differential *B*, *V*, *R* and *V_c*, light curves and, (*B*–*V*) and (*V*–*R*) colour curves of the star HD 81032. Here *V_c* stands for the differential *V* band light

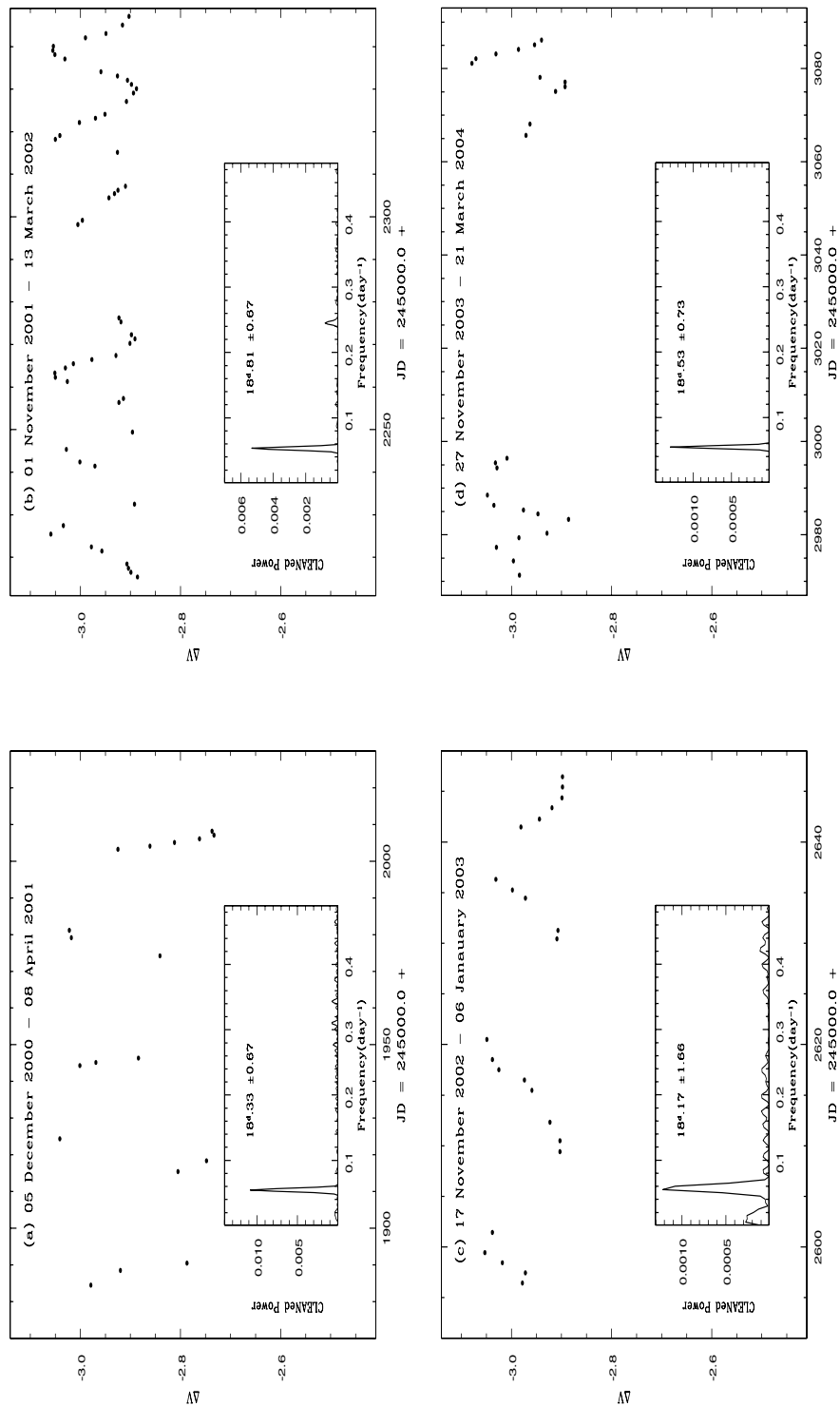


Figure 1. The V band light curves and corresponding CLEANed power density spectra (insets). The epoch is given at the top of each panel, and the period is written at the top of each inset.

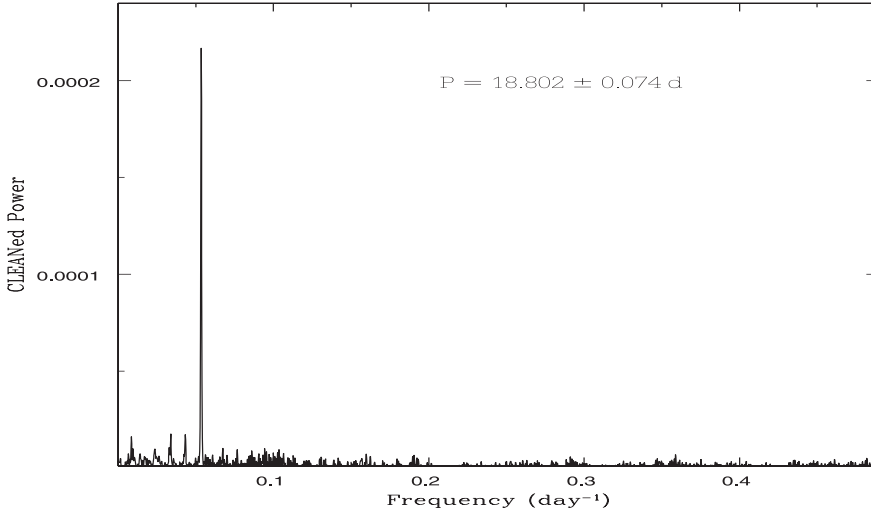


Figure 2. CLEANed power spectra of whole data taken during 2000–2004 of HD 81032.

curve of the comparison and the check stars. Each point in the light curves is mean of 3–4 independent observations taken over a night. The light curve during the observing years 2001–2002 has a dense temporal coverage. We, therefore, divided this light curve into two different epochs to see any variation in the $\theta_{\min}$ and the amplitude. The mean epoch of the light curves, the observed maximum ($\Delta V_{\max}$) and minimum ($\Delta V_{\min}$) in the V band, peak to peak amplitude ($\Delta V = \Delta V_{\max} - \Delta V_{\min}$), and phase of minima ($\theta_{\min}$) are listed in Table 1. The value of $\Delta V_{\max}$ was constant during each epoch indicating that the brightness of unspotted photosphere was constant from epoch to epoch. However, the value of $\Delta V_{\min}$ was reduced by 0.14 mag from epoch ‘a’ to epoch ‘b’, and remained constant during the epoch ‘c’, ‘d’ and ‘e’.

The RS CVn systems usually show one or two well defined minima, thereby indicating that the rotational modulations were caused by one or two prominent spots or groups of spots. Additional spots may be present at other longitudes, or in the circumpolar regions but contribution to the overall rotational modulation may not be appreciable. The phase of the light minimum ($\theta_{\min}$) directly indicates the mean longitude of the dominant groups of spots. A sharp minimum was observed during the epoch ‘a’ (see Fig. 3a). At the same time the amplitude of the V band light curve was found to be 0.288 mag, which was maximum during our observations. The sharpness of the minimum indicate that it is the latitudinal extent of the groups of spots that may be responsible rather than the longitudinal extent. Broad minima during the epochs ‘b’, ‘c’ and ‘d’ indicate that the spots were spread over an appreciable longitudinal range (Fig. 3b to d). It is interesting to see the light curve of the star HD 81032 during the epoch ‘e’. Here a single large spot, characterized by a broad minimum during the epochs ‘b’, ‘c’ and ‘d’ separated into two groups of spots. This can be easily seen by two well separated minima (see Fig. 3e). A significant change in $\theta_{\min}$ (see Table 1) is probably associated with a change in the spot configuration on the surface of the star.

Variation in the colour of the star HD 81032 is correlated with its magnitude (see top two panels of Fig. 3a to e) i.e., the star becomes redder when fainter, and bluer when brighter, supporting the starspot hypothesis. The significance of the correlation has been calculated by determining the linear correlation coefficient, r , between the magnitude and the colours. The value of r between V and $(B-V)$, V and $(V-R)$ and $(B-V)$ and $(V-R)$ was found to be 0.20, 0.55 and 0.40, with the corresponding probability of no correlation being 0.0253 , 6.157×10^{-12} and 2.443×10^{-6} , respectively. The conventional starspot model assumes that the spots are cooler than the surrounding photosphere, and hence one would expect the star to be the reddest at the light minimum, although flaring activity can lead to an opposite effect being observed as seen in UX Ari (Raveendran & Mohin 1995; Padmakar & Pandey 1999; Ulvås & Henry 2003).

Figure 4 shows the plot of the mean epoch versus phase minima ($\theta_{\min}$) of the V band light curves. The following equation was fitted to the data by the method of linear least squares to determine the rate of the phase shift (or migration period).

$$\theta_{\min} = \omega(t - t_0) + \theta_0. \quad (1)$$

Here $\omega = 2\pi/P$ is the rate of phase shift (degrees/day), P is the corresponding period in days, t is time in days, t_0 is reference time, and θ_0 is the reference spot longitude in degree. After the epoch JD 2452400 (or after the epoch 'd'), the evolution of different spots seems to be more responsible than the migration of single group of spots, so, we have used only first three points of the Fig. 4 (i.e., the phase minima of epoch 'a', 'b' and 'c') in the linear least square fit of equation (1). The solid line in Fig. 4 shows the linear least square fit of the equation (1) to the data. The rate of phase shift for the group of spots was determined as 0.1348 ± 0.002 degree/day, which corresponds to a migration period of 7.43 ± 0.07 years. The smaller group of spots during the epoch 'e' appears to rotate with the similar angular velocity with that of a group of spots during the epoch 'a', 'b' and 'c' (see the dotted line in Fig. 4). However, long-term studies of spot groups' evolution and migration in the star HD 81032 require much more frequent sampling in the time domain. The arrangement of groups of spots in one permanent strip can be easily interpreted as a long lived active group rotating with the same velocity. The open circles in the Fig. 4 are not following that strip, and may be due to the formation of a new group of spots on the surface of the star.

5. H α and CaII H and K emission lines

The H α and CaII H and K emission lines are important indicators of chromospheric activity. In very active stars like II Peg, V711 Tau, UX Ari the H α emission is seen clearly above the continuum, whereas in less active stars only a filled in absorption line is observed. Figure 5 shows spectra of HD 81032 in the H α region. Spectrum of the star HD 71952 is shown for comparison. H α emission line can be seen clearly in the star HD 81032 (Fig. 5). It has been suggested that in the RS CVn systems at least one star must also show the intense emission in the H and K lines of CaII (Fekel *et al.* 1986). As shown in Fig. 6, CaII H and K lines are above the continuum at all the phases observed. The measured equivalent widths (EWs) of the H α , CaII H and K emission feature, the corresponding JD and phase are listed in Table 2.

The photometric observations during the epoch 'd' were close to the spectroscopic observations. Figure 7 shows the plot of EWs of CaII H, K and H α against phase,

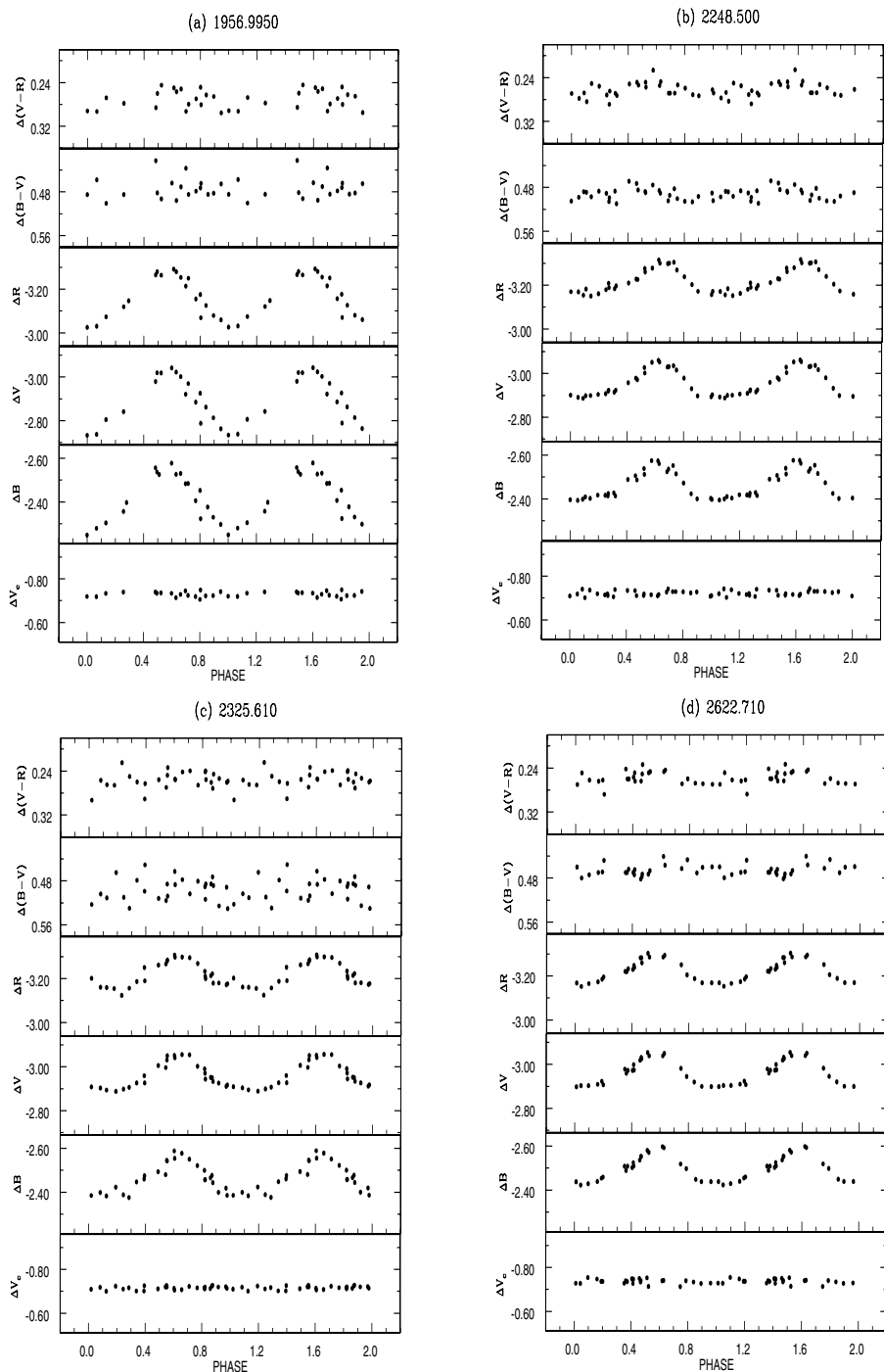
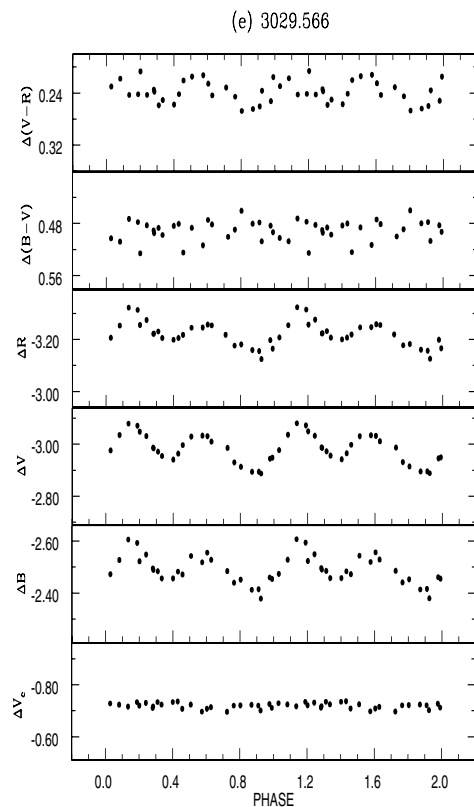


Figure 3. Differential V_c , B , V , R light curve and $(B-V)$, $(V-R)$ colour curve of HD 81032. ΔV_c stands for the differential magnitude between comparison and check stars. The light curve is folded using the period 18.802 d. The epoch of each light curve is mentioned at the top of each figure.

**Figure 3.** (Continued)**Table 1.** Parameters determined from the light curves of the star HD 81032.

Mean epoch 245000.0+	Amplitude	$\Delta V_{\min}$	$\Delta V_{\max}$	Phase minima	
				I	II
(a) 1946.3350	0.288	−2.739	−3.027	0.01	
(b) 2245.8599	0.160	−2.893	−3.053	0.12	
(c) 2322.6527	0.156	−2.896	−3.052	0.15	
(d) 2621.4426	0.152	−2.890	−3.042	0.07	
(e) 2983.8913	0.175	−2.891	−3.066	0.89	0.40

along with the V band light curve of the epoch ‘d’. It appears that $H\alpha$, CaII H , and K emission features are variable and anti-correlated with the photometric phase, i.e., the maximum at photometric minimum and minimum at photometric maximum. This could be due to the presence of active cool spots on the surface of the star.

6. Spectral type

The value of the total galactic reddening $E(B-V)$ in the direction of HD 81032 is estimated to be 0.04 mag from the reddening map given by Schlegel *et al.* (1998).

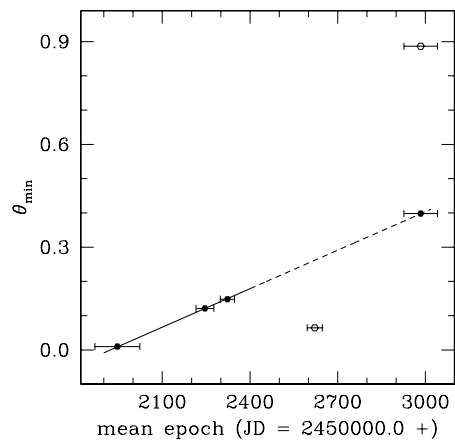


Figure 4. Plot of mean epoch versus light minima. The solid line represents the linear least square fit of the equation 1. The open circles were not used in the fit. The horizontal bar represents the length of the epoch.

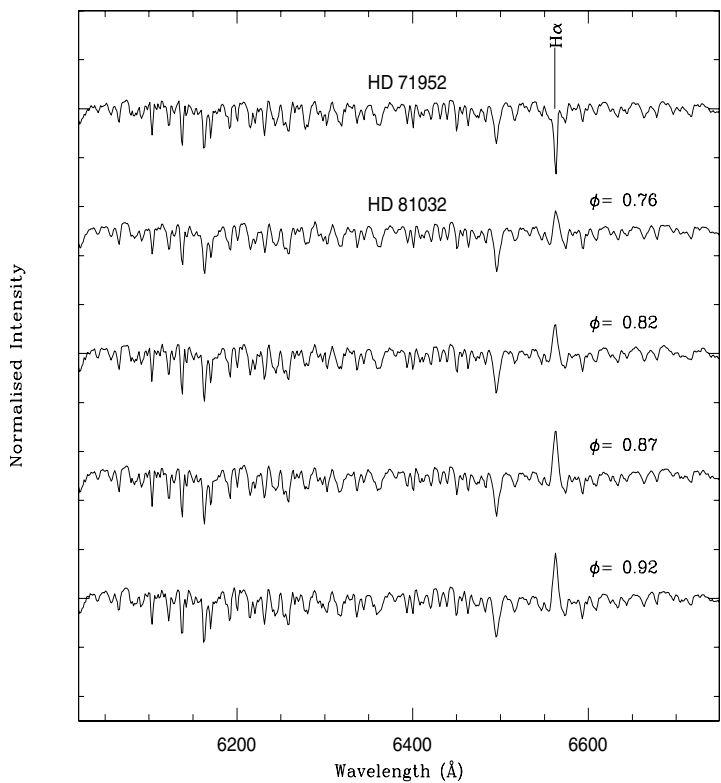


Figure 5. H α spectra of HD 81032. HD 71952 was taken as reference spectra.

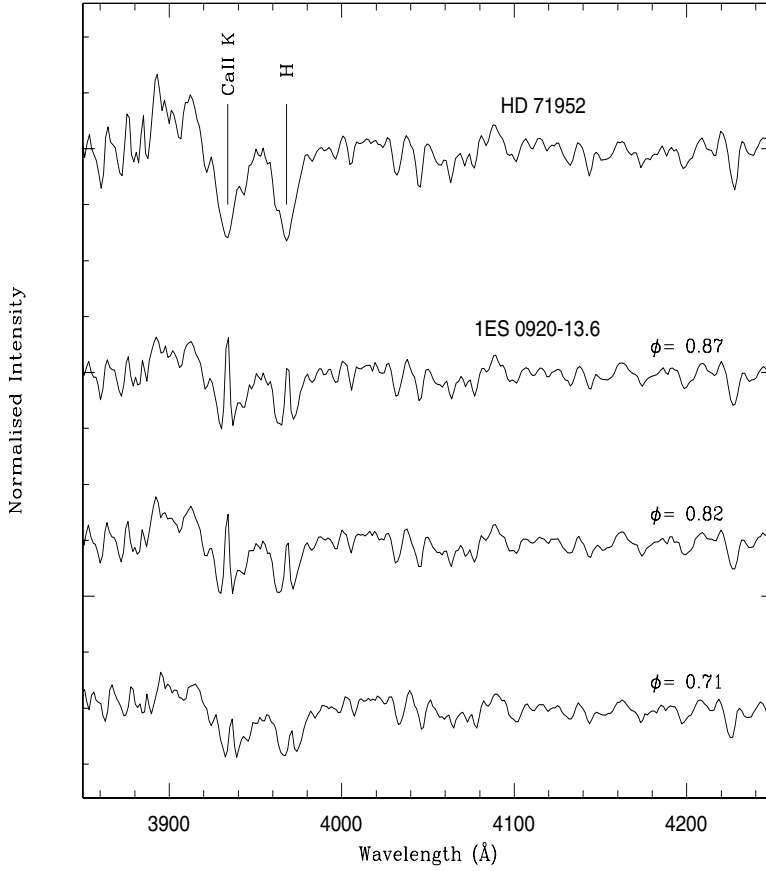


Figure 6. CaII H and K spectra of HD 81032. Spectra of HD 71952 was taken as a reference.

Table 2. Equivalent widths (EWs) of CaII H and K and H α emission lines.

JD	Phase	EWs (Å)		
		CaII K	H	H α
245000+				
2659.626	0.71	1.828	1.037	–
2660.513	0.76	–	–	0.980
2661.520	0.82	4.393	2.661	1.178
2662.518	0.87	5.012	2.949	1.651
2663.536	0.92	–	–	1.680

However, the star HD 81032 suffers negligible reddening at a distance of 140 pc, and we assume $E(B-V) = 0$ for the remaining discussion. Using the Tycho parallax (Hog 1997) and the value of $V (= 8.63 \text{ mag})$ mag from the present photometry the absolute magnitude M_V of HD 81032 is 3.04 mag. The value of M_V is consistent with the luminosity class IV for this star. The $(B-V)$ colour of HD 81032 derived from

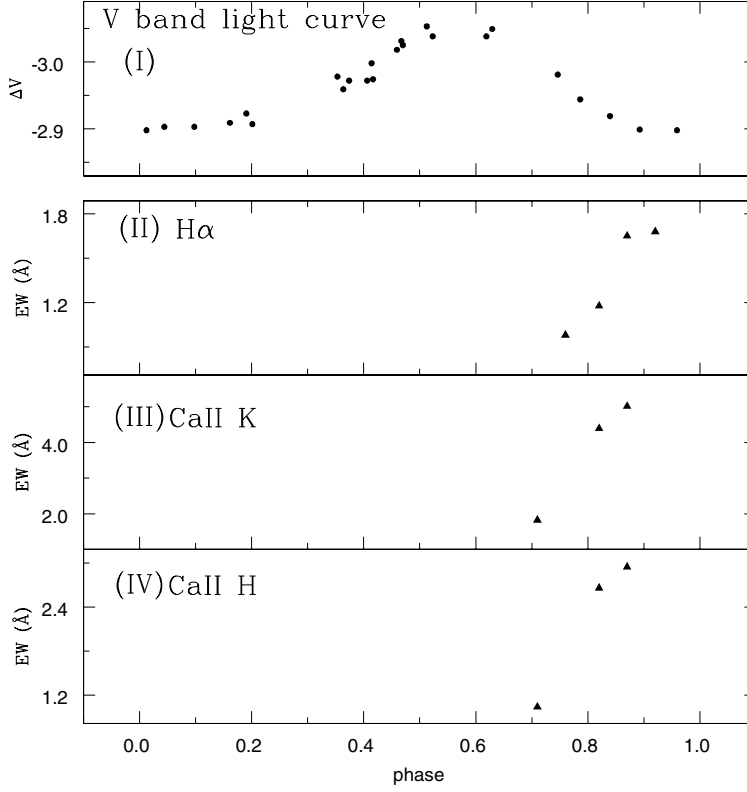


Figure 7. (I) V light curve of the HD 81032 during the epoch ‘d’, (II) variation of EWs of H α emission feature, (III) CaII K EWs and (IV) CaII H EWs. Phases are reckoned from JD 2452307.761 and the period $18^d.802$.

our photometry is 1.02 ± 0.01 consistent with the relatively precise to the value of 0.98 ± 0.05 given in the Tycho catalogue (Wright *et al.* 2003). The $(B-V)$ colour is best matched with the spectral class K0 IV, and is consistent with the spectral class identified by Wright *et al.* (2003). Applying the bolometric correction -0.4 (Schmidt-Kaler 1982) for K0 IV star, the bolometric magnitude of HD 81032 is 2.65 mag.

7. Spectral energy distribution (SED)

We have determined the SED of HD 81032 using broad band UBVRI (present photometry), 2MASS JHK (Cutri *et al.* 2003) and 12, 25, 60 and 100μ IRAS (Moshir 1989) fluxes. However, only upper limits are available at 25, 60 and $100 \mu\text{m}$. The observed SED of HD 81032 along with the synthetic SED is shown in Fig. 8. The synthetic SED is expected from the intrinsic properties of the star (Kurucz 1993). The model SED shown in Fig. 8 has been adjusted to coincide with the observed SED of the star HD 81032 at the V-band wavelength of 0.55μ . We have overplotted the synthetic SED for the different T_{eff} and $\log g$ combinations. The values of T_{eff} and $\log g$ which best match the observed SED are $5000 \pm 250 \text{ K}$ and 3.5 ± 0.5 , respectively, and are consistent within one subclass of the inferred K0 IV spectral type of the star HD 81032 (see section 6.).

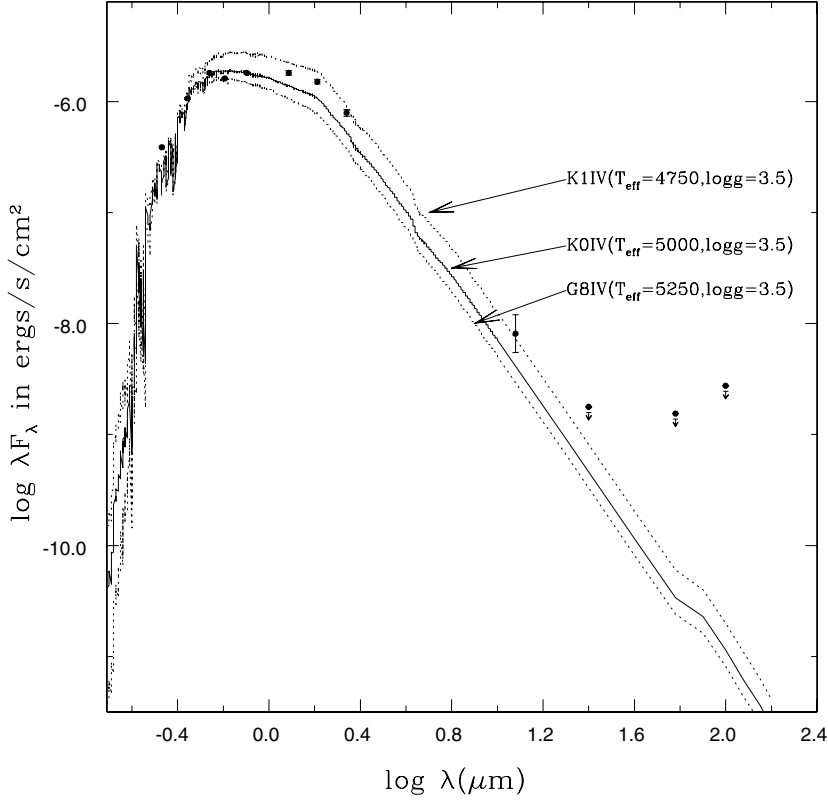


Figure 8. SED of the star HD 81032 (solid dots). The solid line represents the best match model SED from Kurucz (1993) as expected from the intrinsic properties of the star, dotted lines represent the SED of the K1IV and G8IV type star.

The continuum of HD 81032 is in good agreement with the normal values for the spectral type of K0 IV up to JHK band but deviates in all bands longer than $25 \mu\text{m}$ (see Fig. 8). The difference between the observed and the expected (J–K) and (H–K) colours are 0.06 ± 0.04 and 0.005 ± 0.03 mag, respectively. This indicates that there is no significant colour excess in J, H and K band. The expected (J–K) and (H–K) colours were taken from the spectral type of the star (Koorneef 1983). The $12 \mu\text{m}$ magnitude (m_{12}) of the star HD 81032 was found to be 6.09 ± 0.17 mag using the relation $m_{12} = -2.5 \log f_{12} + 3.63$ (Mitrou *et al.* 1996). The intrinsic $[K - 12]$ colour of the star HD 81032 is determined to be 0.03 ± 0.17 mag. The expected $[K - 12]$ colour of 0.14 mag for the K0 IV type star (Verma *et al.* 1987) gives the colour excess of 0.17 mag, which is consistent with the ideal value of 0.0 ± 0.2 mag (Mitrou *et al.* 1996).

8. X-ray light curve and spectra

The light curve for the source and background were extracted using the *xselect* package for the PSPC 0.1–2.4 keV energy band which contains all the X-ray photons. The

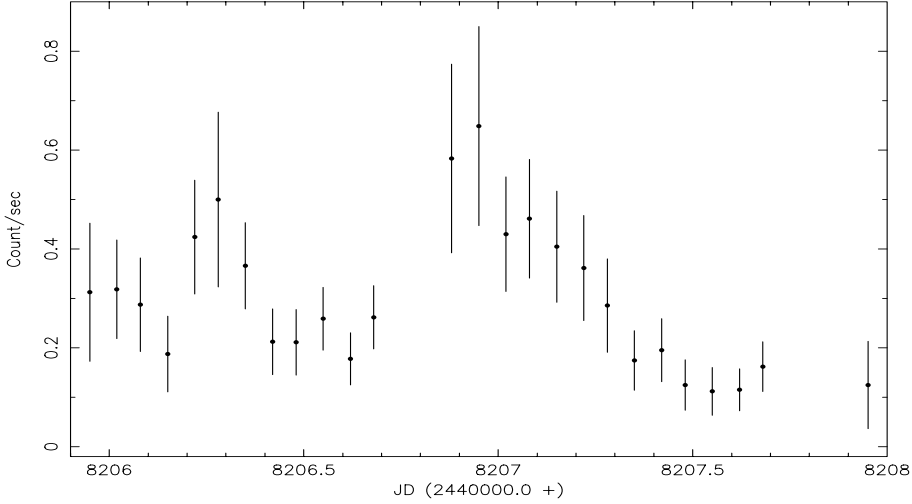


Figure 9. Background subtracted RASS PSPC X-ray light curve of HD 81032. The bin size in the light curve is 64 s.

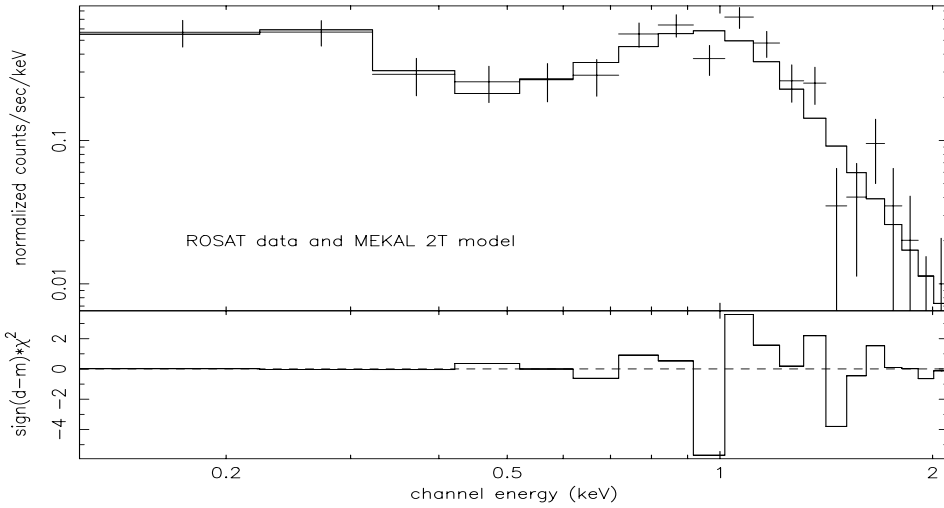


Figure 10. Spectra of HD 81032 taken with the ROSAT PSPC detector (top panel), the bottom panel shows the χ^2 for each bin.

background-subtracted X-ray light curve of HD 81032 is shown in Fig. 9 plotted with a time bin size of 64 s. It appears from the light curve that a moderate flare occurred during the RASS observations, with a peak of about 0.6 ct s^{-1} (compared to a pre-flare level of $0.2\text{--}0.3 \text{ ct s}^{-1}$) at approximately $\text{JD} = 2448206.95$ (1990/11/11 10:48:0.0) and a half-decay time of $2.6 \times 10^4 \text{ s}$.

The X-ray spectrum of HD 81032 as observed with the ROSAT PSPC is shown in Fig. 10. Response matrices based on the available off-axis calibration of the PSPC and using the appropriate ancillary response files were created and data were fitted using the *xspec* (version 11.3.1) spectral analysis package. Spectral models for

Table 3. Results of X-ray spectral analysis.

Model	Abundances ^a	N_H 10^{20} cm^{-2}	kT_1 (keV)	EM_1 10^{52} cm^{-3}	kT_2 (keV)	EM_2 10^{53} cm^{-3}	χ^2_ν	Degrees of freedom
MEKAL 1T	1.0 (fixed)	$0.21^{+0.22}_{-0.16}$	$0.85^{+0.29}_{-0.10}$	$5.2^{+2.2}_{-2.5}$			1.87	17
MEKAL 1T	$0.19^{+0.14}_{-0.08}$	$1.5^{+1.0}_{-1.0}$	$0.84^{+0.20}_{-0.17}$	$18.9^{+14.7}_{-9.9}$			1.53	16
MEKAL 2T	1.0 (fixed)	$1.4^{+5.4}_{-1.1}$	$0.2^{+0.2}_{-0.1}$	$2.9^{+5.7}_{-1.5}$	$1.12^{+0.52}_{-0.36}$	$1.8^{+1.1}_{-1.6}$	1.4	15

^aCommon value of abundances for all the elements respect to the solar photosphere values; errors are with 90% confidence for single parameter based on $\chi^2_{\min} + 2.71$; distance = 140 pc.

thermal equilibrium plasma known as the Mewe–Kaastra–Liedahl or MEKAL model (Liedahl *et al.* 1995; Mewe *et al.* 1995) were used. The background-subtracted X-ray spectra were fitted with 1-temperature (1T) and 2-temperature (2T) plasma models, either assuming solar photospheric abundances as given by Anders & Greves (1989) or allowing the abundance of every element other than H to vary by a common factor relative to the solar (photospheric) values. In each of the above models the interstellar absorption was assumed to follow the absorption cross-sections given by Morrison & McCammon (1983), and the total intervening hydrogen column density N_H was allowed to vary freely.

The results of different model fits are summarized in Table 3. Single-temperature MEKAL models with abundances fixed to the solar values gave unacceptably high values for χ^2_ν , and thus can be rejected. However, single-temperature plasma models with abundances of $0.19^{+0.14}_{-0.08}$ times solar and plasma temperature of $0.84^{+0.17}_{-0.20}$ were found acceptable. Alternatively, two-temperature MEKAL models with fixed solar abundances were also found to be acceptable; the two temperatures being $0.2^{+0.2}_{-0.1}$ keV and $1.12^{+0.52}_{-0.36}$ keV. Multi-temperature variable abundance models would also produce acceptable fits to the PSPC spectrum, but the small number of accumulated counts does not warrant such complex models. The PSPC spectrum and the best-fit two-temperature plasma model with solar abundances are shown in Fig. 10 along with the significance of the residuals in terms of their χ^2 .

Based on the best fit 2T MEKAL model a source flux of $9.1^{+0.9}_{-1.6} \times 10^{-12}$ erg cm $^{-2}$ s $^{-1}$ was obtained. At a distance of 140 pc, the X-ray luminosity of HD 81032 is calculated to be $2.1^{+0.2}_{-0.4} \times 10^{31}$ ergs $^{-1}$, similar to that of known subgiants RS CVn binaries (Drake *et al.* 1989; Singh *et al.* 1995, 1996; Dempsey *et al.* 1997; Padmakar *et al.* 2000; Pandey *et al.* 2005). The derived X-ray luminosity is about a factor of 2 bigger than the earlier estimate given in section 1. This may be due to the non-trivial column density of 1.4×10^{20} cm $^{-2}$ inferred: objects with a high column density will typically have a higher count rate to flux conversion factor than the standard one which assumes a low column density. The values of kT_1 and kT_2 of HD 81032 (see Table 3) are consistent with those of other RS CVn systems studied by Dempsey *et al.* (1997), who found an average value of 0.18 ± 0.01 keV and 1.37 ± 0.04 keV for a sample of 28 RS CVn systems. The volume emission measures EM1 and EM2 of HD 81032 are also found to be consistent with those of the other similar systems, where the average volume emission measures EM_1 and EM_2 are $1.2 \pm 0.3 \times 10^{53}$ and $5.7 \pm 1.2 \times 10^{53}$ cm $^{-3}$, respectively (Dempsey *et al.* 1997).

All the errors quoted in this section are internal statistical errors, and that the distance-dependent errors of 45% in luminosity which were mentioned in section 1 have not been included in, for example, the EM errors quoted in Table 3 or the errors in the X-ray luminosity given in the above paragraph of this section.

9. Summary

A photometric period of 18.802 ± 0.074 d has been discovered in the star HD 81032. The amplitude and the phase of the photometric light curves of HD 81032 are observed to be changing from one epoch to another. The amplitude of variation is found to be maximum (0.29 mag) during the observing year 2000–2001. The change in the amplitude is mainly due to a change in the minimum of the light curve, and this may be

due to a change in the spot coverage on the surface of the star. A single group of spots is found to migrate, and a migration period of 7.43 ± 0.07 years is determined from the first three years of the data. The star is seen to be redder at the light minimum and we interpret that this is due to the relatively cooler temperature of the darker regions present in the visible hemisphere. A spectral type of K0 IV is determined for HD 81032.

The equivalent widths of H α , CaII H and K emission lines in the spectra of HD 81032 appear more intense close to the photometric light minimum. This could be attributed to the presence of cool spots on the surface of the stars. The observed X-ray spectrum and the inferred coronal plasma parameters for HD 81032 are typical of those seen in active stars such as RS CVn binaries.

All of the optical and X-ray properties found for HD 81032 are most consistent with it being an evolved RS CVn binary of the long-period type. Although there are a handful of apparently single evolved stars, e.g., FK Com, which can exhibit similar properties, identification of HD 81032 as such a rare object seems rather improbable. Clearly, high-resolution optical spectroscopy of HD 81032 in order to confirm the presence of radial velocity variability and determine a spectroscopic binary orbit would be definitive evidence for its proposed binary nature.

Acknowledgements

We are grateful to the anonymous referee for valuable comments and suggestions. We are thankful to the time allocation committee for giving time at 234-cm VBT. This research has made use of the data obtained from the High Energy Astrophysics Science Archive Research Center (HEASARC), provided by NASA's Godard Space Flight Center. Starlink is funded by PPARC and based at the Rutherford Appleton Laboratory, which is part of Council for the Central Laboratory of the Research Councils, UK. This publication makes use of data products from the Two Micron All Sky Survey, which is a joint project of the University of Massachusetts and the Infrared Processing and Analysis Center/California Institute of Technology, funded by the National Aeronautics and Space Administration and the National Science Foundation.

References

- Anders, E., Grevesse, N. 1989, *Geochim. Cosmochim. Acta*, **53**, 197.
- Ayres, T. R. *et al.* 1995, *ApJS*, **96**, 223.
- Cutri, R. M. *et al.* 2001, 2MASS All-Sky Catalog of Point Sources, VizieR On-line Data Catalog II/246.
- Dempsey, R. C., Linsky, J. L., Fleming, T. A., Schmitt, J. H. M. M. 1997, *AJ*, **478**, 358.
- Drake, S. A., private communication.
- Drake, S. A., Simon, J., Linsky, J. L. 1992, *ApJS*, **82**, 31.
- Drake, S. A., Simon, J., Linsky, J. L. 1989, *ApJS*, **71**, 905.
- Drake, S. A., Linsky, J. L. 1986, *AJ*, **91**, 1229.
- Elvis, E., Plummer, D., Schachter, J., Fabbiano, G. 1992, *ApJS*, **80**, 257.
- Fekel, F. C., Moffet, T. J., Henry, G. W. 1986, *ApJS*, **60**, 551.
- Güdel, M. 2002, *ARAA*, **40**, 217.
- Hog, E. *et al.* 1997, *A&A*, **323**, L57.
- Houk, N., Smith-Moore, M. 1988, Michigan Catalogue of Two-Dimensional Spectral Types for the HD Stars, Vol. 4: -26 to -12 Degrees.
- Koorneef, J. 1983, *A&A*, **128**, 84.
- Kurucz, R. L. 1993, private communication.

- Landolt, A. U. 1992, *AJ*, **104**, 340.
- Liedahl, D. A., Osterheld, A. L., Goldstein, W. H. 1995, *ApJ*, **438**, L115.
- Mewe, R., Kaastra, J. S., Liedahl, D. A. 1995, *Legacy*, **6**, 16.
- Mitrou, C. K., Doyle, J. G., Mathioudakis, M., Antonopolou 1996, *A&AS*, **115**, 61.
- Morrison, R., McCammon, D. 1983, *ApJ*, **270**, 119.
- Moshir, M. 1989, Explanatory Supplement to IRAS Faint Source Survey.
- Padmakar, P. S., Singh, K. P., Drake, S. A., Pandey, S. K. 2000, *MNRAS*, **314**, 733.
- Padmakar, P. S., Pandey, S. K. 1999, *A&AS*, **138**, 203.
- Pandey, J. C., Singh, K. P., Drake, S. A., Sagar, R. 2005, *AJ*, **130**, 1231.
- Pandey, J. C., Singh, K. P., Sagar, R., Drake, S. A. 2002, *JApA*, **23**, 9.
- Pfeffermann, E. et al. 1987, In: *Soft X-ray Optics and Technology, Proc. SPIE*, **733**, 519.
- Raveendran, A. V., Mohin, S. 1995, *A&A*, **301**, 788.
- Roberts, D. H., Lehar, J., Dreher, J. W. 1987, *AJ*, **93**, 968.
- Schachter, J. F. et al. 1996, *ApJ*, **463**, 747.
- Schlegel, D. J., Finkbeiner, D. P., Davis, M. 1998, *ApJ*, **500**, 525.
- Schmidt-Kaler Th., 1982, In: Landolt-Bornstein, Numerical Data and Functional Relationship in Science and Technology, New series, Group VI, Vol. 2b, (eds) Scaifers K., Voigt H. H., Springer-Verlag, Berlin.
- Schmitt, J. H. M. M., Liefke, C. 2004, *A&A*, **417**, 651.
- Singh, K. P., Drake, S. A., White, N. E. 1995, *ApJ*, **445**, 840.
- Singh, K. P., Drake, S. A., White, N. E., Simon, T. 1996, *AJ*, **112**, 221.
- Trümper, R. A. 1983, *Adv. Space Res.*, **2**, 241.
- Ulvås V. Aarum, Henry, G. W. 2003, *A&A*, **402**, 1033.
- Verma, R. P., Iyengar, K. V. K., Rengarajan, T. N. 1987, *A&A*, **177**, 346.
- Voges, W. et al., 1999, *A&A*, **349**, 389.
- Wright, C. O., Egan, M. P., Kraemer, K. E., Price, S. D. 2003, *AJ*, **125**, 359; *A&A*, **172**, 225.

Spectral Analysis of the Background in Ground-based, Long-slit Spectroscopy

Frédéric Zagury

Institut Louis de Broglie, 23 rue Marsoulan, 75012 Paris, France.

e-mail: fzagury@wanadoo.fr

Received 2004 September 27; accepted 2005 August 24

Abstract. This paper examines the variations, because of atmospheric extinction, of broad-band visible spectra, obtained from long-slit spectroscopy, in the vicinity of some stars, nebulae, and one faint galaxy.

Key words. Atmospheric effects—planetary nebulae: general—planetary nebulae: individual (HD 44179, NGC 2022, NGC 6309, NGC 6891).

1. Introduction

In a preceding paper (Zagury & Goutail 2003) we analyzed the extinction effects of the atmosphere on sunlight, as they can be observed by a balloon-borne spectrometer. The study specified the radiation field, due to the Sun, in any point of a clear atmosphere, and we were able to derive its analytical form.

The effect of the atmosphere on the visible continuum of an external source of light is three-fold: Rayleigh extinction by the gas (nitrogen mainly, $\propto e^{-g/\lambda^4}$); extinction by aerosols ($\propto e^{-a/\lambda}$); and the broad-band ozone absorption (Chappuis bands).

In the complete forward direction, direct light from the source will be extinguished according to these three effects. In near-forward directions, direct light from the source is still present due to refraction and turbulence in the atmosphere, but, as we move away from the direction of the source, it drops in intensity, and the source's light scattered by aerosols (with a spectrum $\propto e^{-a/\lambda}/\lambda$) will become discernible. Farther away, when direct light from the source is negligible, forward scattering by aerosols and ozone absorption will be the only remaining processes at work. At very large angles, forward scattering is no more effective; isotropic Rayleigh scattering by the gas ($\propto e^{-g/\lambda^4}/\lambda^4$) and the ozone absorption feature should dominate the spectrum. A first application of this work was presented by Zagury & Fujii (2003) who were able to fit the spectra of a red horizon at sunrise.

In this paper, I shall be concerned with the continuum of spectra observed in the vicinity of different types of objects: stars, nebulae and galaxies. The spectra, obtained through long-slit spectroscopy, are considered after correction for the instrumental response, but before absolute calibration (no correction of atmospheric extinction), so that the effects mentioned above can be recognized and eventually separated.

Sections 2. and 3. present the different observations used in the paper and a series of notes collected during the investigations for its preparation. Observations are analyzed in sections 4. and 5. and discussed in section 6..

2. Observations

The observations used in this paper all emanate from long slit spectroscopy. Some were observed with the FAST spectrograph (Fabricant *et al.* 1998) using the 1.5 m Tillinghast telescope of the Fred Lawrence Whipple Observatory (Mount Hopkins, U.S.A). The others come from the Kwitter–Henry database (<http://cf.williams.edu/public/nebulae/browse.cfm>; Kwitter & Henry 2002), and were observed with the 2.1 m telescope at Kitt Peak National Observatory (KPNO). I will also refer to night sky spectra observed by Phil Massey and Craig B. Foltz (Massey & Foltz 2000), also obtained at Kitt Peak and Mount Hopkins.

Table 1 gives the positions and the visual magnitudes of the objects (central stars for nebulae) used in the paper. Spectrum ‘x’ of a 2-D array will be noted *sx*. Vertical dotted lines on the plots give the limits of the ozone absorption region.

2.1 Mount Hopkins FAST observations

The reference paper for the FAST spectrograph is Fabricant *et al.* (1998). The configuration used for the observations is a 300 line mm^{-1} grating. The slit samples a region $\sim 3'$ long and $3''$ wide on the sky. The wavelength coverage goes from 3660 Å to 7530 Å, with a resolution of 1.47 Å/pixel . The slit is divided into 320 pixels, with a spatial resolution of $0.61''/\text{pixel}$. Note that the level of undesired scattered light (reflections from the refractive optics and the CCD, imperfections in the reflective surfaces and scattering of light outside first order from the spectrograph case) is estimated to be less than 1% of the spectra (see the FAST specifications at <http://tdc-www.harvard.edu/instruments/fast/>).

The observations (Table 2) of the three stars (HD 23302 (17 τ), HD 44113, HD 44179) and of the Red Rectangle nebula are extracted from a larger set observed by Dr. Lucas Macri during three observing runs, on December 21 and 22, 2001; February 9, 2002 and March 26, 2003. Observing conditions during the three runs were good, with a seeing in the range $1\text{--}2''$. Galaxy UGC 11917 was observed by Dr. I. Ginsburg on October 31, 2002.

Table 1. Coordinates, visual magnitudes, size of objects used in the paper.

Object	α_{2000}	δ_{2000}	m_V	size ⁽¹⁾
HD 23302	03 44 52	+24 06 48	3.7	
HD 44113	06 19 37	−11 03 08	~ 9	
HD 44179	06 19 58	−10 38 15	9.0	26 ⁽²⁾
NGC 2022	05 42 06	+09 05 10	10.1	19 ⁽³⁾
NGC 6309	17 14 04	−12 54 36	11.6	15.5 ⁽³⁾
NGC 6891	20 15 09	+12 42 17	12.4	15 ⁽³⁾
UGC 11917	22 08 08	+45 34 15	~ 17	24 ⁽⁴⁾

⁽¹⁾ Size of extended objects ($''$).

⁽²⁾ Size of the Red Rectangle nebula $11''$ from HD 44179.

⁽³⁾ Size of nebula, from Acker *et al.* (1992).

⁽⁴⁾ Minor axis, in the red, from the Uppsala Galaxy Catalogue (UGC).

Table 2. Mount Hopkins FAST observations.

Position	U.T. ⁽¹⁾	Δt ⁽²⁾	A.M. ⁽³⁾	Alt. ⁽⁴⁾
December 2001 (2001-12-21)				
HD 23302-1	06:08:49	0.2	1.04	87.18
HD 23302-2	06:13:10	0.2	1.04	87.32
December 2001 (2001-12-22)				
HD 44179-1	08:44:21	15	1.41	44.92
R.R. nebula ⁽⁵⁾	09:03:31	120	1.46	42.96
HD 44179-2	09:10:04	15	1.48	42.44
February 2002 (2002-02-09)				
HD 44113	5:05:06	15	1.38	46.28
HD 44179	5:07:24	5	1.37	46.59
March 2003 (2003-03-26)				
HD 23302	03:02:00	0.2	1.71	35.84
HD 44179-1	3:10:46	5	1.51	41.40
HD 44179-2	3:33:27	5	1.60	38.52
October 2002 (2002-10-31)				
UGC 11917	04:59:03	1200	1.17	58.89

(1) UT HH:MM:SS at start of exposure.

(2) Duration of exposure (seconds).

(3) Air mass.

(4) Altitude ($^{\circ}$) of the object at time of observation.

(5) Observed 11.5' north from HD 44179.

The data reduction was done by Dr. Susan Tokarz (Tokarz & Roll 1997). Wavelength calibration is made with a HeNeAr lamp. A final 2-D array is extracted for each observation, which is dark-corrected and corrected for the instrumental response, but not for atmospheric extinction. The 2-D arrays have been binned by two or by four during the process, along the spatial dimension, yielding a final spatial resolution of 1.2''/pixel for the binned by two spectra and 2.4''/pixel for the binned by 4. All 2-D FAST spectra, except those from February 2002, are binned by two. 1-D spectra of point source objects are usually extracted from the 2-D arrays (section 3.2).

2.2 Kitt Peak observations

The Kwitter–Henry database (<http://oit.williams.edu/nebulae/browse.cfm>) proposes a collection of 88 nebular spectra in the [3600 Å, 9600 Å] wavelength range. Dr. Kwitter gave me access to the 2-D arrays, at the same early stage of data reduction as for the Mount Hopkins observations (i.e., corrected for the instrumental response but not for atmospheric extinction), of three planetary nebulae: NGC 2022, NGC 6309, and NGC 6891.

These spectra, as for the preceding FAST spectra, are issued from long-slit spectroscopy (see Kwitter & Henry 2001). The slit is 5'' wide, 285'' long in the E–W direction. The spatial resolution is 0.78''/pixel. The spectral resolution is 1.49 Å/pixel. Only the blue spectra will be used, which range from 3550 Å to 6800 Å for the NGC 2022 observation, from 3640 Å to 6780 Å for NGC 6309 and NGC 6891 observations. Contrary to the FAST spectra, these spectra are unbinned.

Table 3. Kitt Peak observations.

Object	Date ⁽¹⁾	U.T. ⁽¹⁾	Δt ⁽²⁾	$\Delta \delta$ ⁽³⁾	A.M. ⁽⁴⁾	Alt. ⁽⁵⁾
NGC 2022	1996-12-08	06:52:31	300	$-6.5^{(6)}$	1.12	62.41
NGC 6309	1999-07-01	04:21:08	300	$\sim +3$	1.59	38.67
NGC 6891	1999-07-01	09:34:58	30	~ 1	1.07	69.60

⁽¹⁾ UT date (YYYY-MM-DD) and time (HH:MM:SS) at the start of the exposure.

⁽²⁾ Duration of exposure (seconds).

⁽³⁾ Declination offset from the central star (").

⁽⁴⁾ Air mass.

⁽⁵⁾ Altitude (°) of the object at the time of observation.

⁽⁶⁾ Requested offset was $6.5''$ S. The header, which normally indicates the central position of the slit, gives $16''$ S.

The observations were obtained at KPNO with the 2.1 m telescope, July 1, 1999 for NGC 6309 and NGC 6891 (Kwitter & Henry 2001), December 8, 1996 for NGC 2022 (Kwitter *et al.* 2003) (Table 3).

3. Data overview

3.1 2-D spectra of stars

Figure 1 plots spectra from the 2-D array, after instrumental calibration and before correction for atmospheric extinction, of one of the stars (HD 44179, first March 2003 observation). Although the direct light from HD 44179 is mainly concentrated in one or two pixel(s) (pixels 52 and 51 in this example), which corresponds to the seeing, and fixes the direction of the star, it is distributed over several pixels (here pixels 49 to 54, which represent $\sim 7''$ on the sky), with a decrease in intensity of a factor of 100 between pixel 52 and pixels 49 or 54.

The main cause of the spread of the signal over several central pixels is the dispersion of light in the atmosphere due to turbulence and refraction. The effect of atmospheric extinction on these spectra is manifest in the blue, the long wavenumber (short wavelength), $1/\lambda > 2 \mu\text{m}^{-1}$, bending of the spectra being due to Rayleigh extinction by nitrogen ($\propto e^{-g/\lambda^4}$).

Ozone absorption affects the $[1.5 \mu\text{m}^{-1}, 2 \mu\text{m}^{-1}]$ wavenumber region (Fig. 2), but is not always easy to distinguish. Modifications of the sun spectrum in Zagury & Goutail (2003), due to ozone, are perceptible for column densities of $\sim 9 \cdot 10^{18}$ molecule/cm², and are important when the column density reaches a few 10^{19} molecule/cm². Zenithal ozone column densities observed from Earth are of order 300 Dobson (1 Dobson = $2.69 \cdot 10^{16}$ molecule/cm²). For zenithal distances of order $\theta \sim 45^\circ$, as for HD 44113 and HD 44179 observations, the ozone column density will be $\sim 8 \cdot 10^{18} / \cos \theta = 10^{19}$ molecule/cm². Hence, our observations are in the transition region where ozone absorption begins to be significant.

The differences between the main spectra of an observation (pixels 49 to 55 in Fig. 1) can be appreciable, and there is an order in the successive modifications of the spectrum's shape. Moving from spectrum *s*49 to *s*54, there is a progressive decrease (in absolute value) of the average slope of the spectra; this coincides with, and may

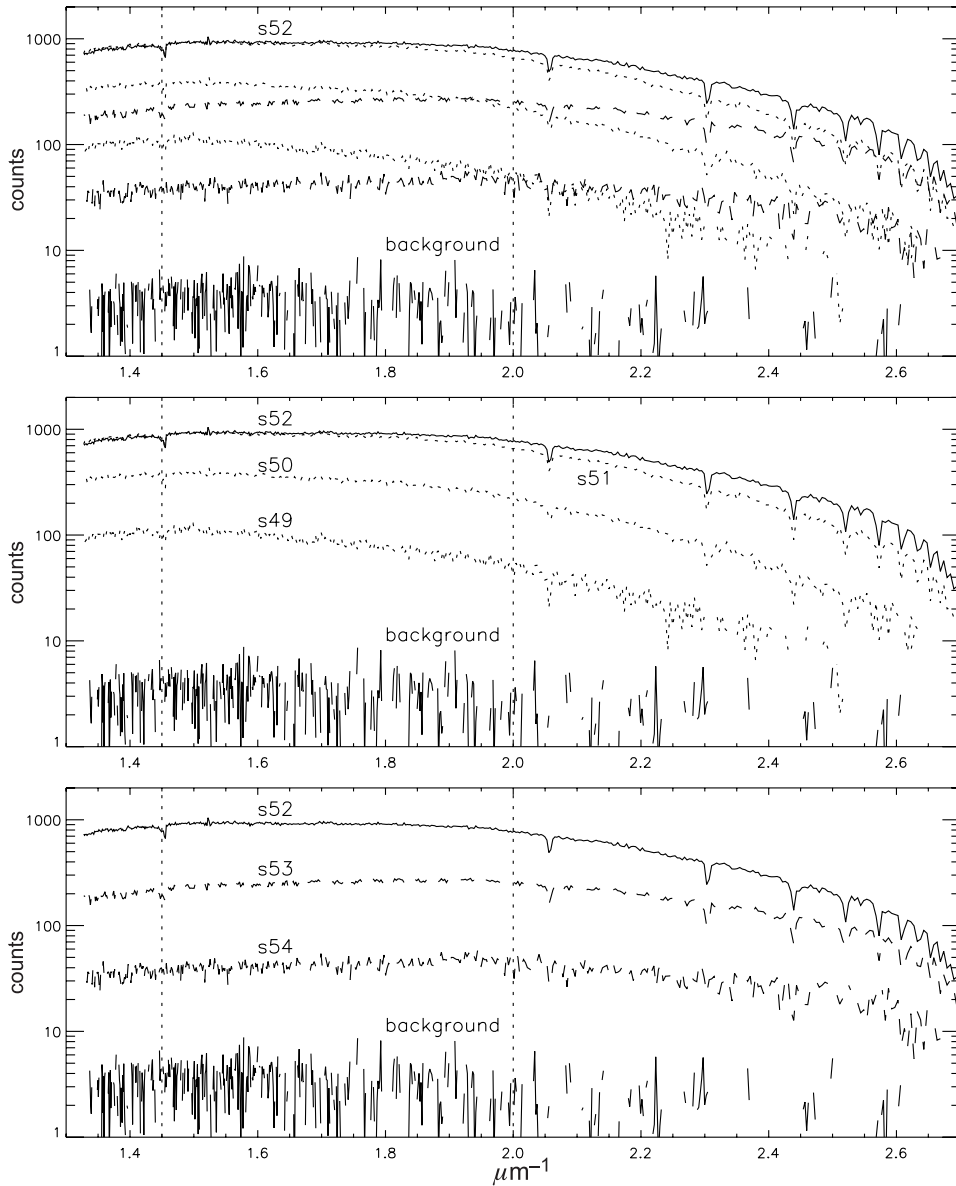


Figure 1. HD 44179 (Red Rectangle), first March 2003 observation. The pixel with maximum signal (main pixel of the 2-D array) is pixel 52. Most of the direct starlight is concentrated in pixels 49 to 54. The spectra then reach a common shape (the background).

be due to, the orientation of the slit in the azimuth-elevation plane (pixel numbers increase with pixel altitudes). This is observed in all star observations.

The relationship between these spectra often takes a simple analytical expression. Main spectra in Fig. 1, for instance, relate to each other by an exponential of $1/\lambda^4$ (as for the second March 2003 observation of HD 44179, Fig. 3) which could be due to the differences in extinction by the gas (though the coefficients of the exponent are

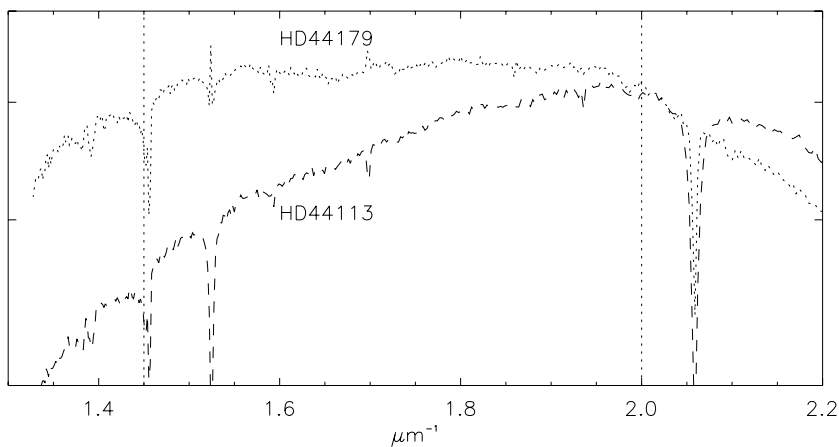


Figure 2. 1-D spectra of February 2002 observations of HD 44179 and HD 44113 (arbitrary scaling). Ozone absorption is responsible for the flattening of HD 44179 spectrum and the depression between $\sim 1.5 \mu\text{m}^{-1}$ and $\sim 1.9 \mu\text{m}^{-1}$. Although present in the same proportion, it is more difficult to detect in the observation of HD 44113 because of the steep rise of the spectrum. Rayleigh $1/\lambda^4$ extinction is responsible for the decrease longward of $2 \mu\text{m}^{-1}$.

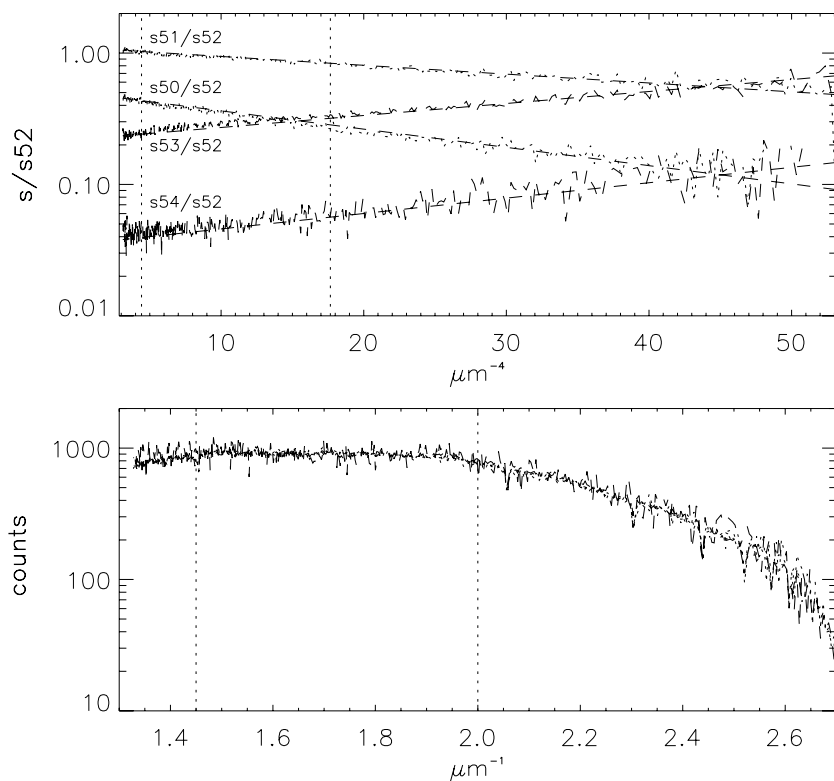


Figure 3. Relations between the spectra in the second March 2003 observation of HD 44179. Top: ratios of s_{51} , s_{50} , s_{53} , s_{54} to s_{52} are an exponential of $1/\lambda^4$. Bottom: s_{52} is superimposed by $0.9e^{0.015\lambda^{-4}}s_{51}$, $2e^{0.032\lambda^{-4}}e^{10^{19}\sigma}s_{50}$, $4.5e^{-0.021\lambda^{-4}}e^{-2 \cdot 10^{19}\sigma}s_{53}$, and $28.6e^{-0.027\lambda^{-4}}e^{-2 \cdot 10^{19}\sigma}s_{54}$.

high compared to what was found in Zagury & Goutail 2003) and a small correction for absorption by ozone. The transformations hold good for other observations made during the same night and in the same region.

The angular extent over which direct light from a star can be detected depends on the observation. Its diameter, centered on the direction of the star, is of $\sim 12''$ and $\sim 14''$ for the HD 23302 and HD 44179 December 2001 observations. For the March 2003 observation, the extent keeps the same order of magnitude for HD 23302, but as mentioned above, it is divided by 2 for HD 44179. February observation of HD 44179 gives, as for March 2003, $\sim 7''$. These important differences, for HD 44179, between December 2001 and the other runs, could be related to the exposure time (February 2001 and March 2003 exposures are one third of December 2001's), maybe an effect of the time-dependence of turbulence.

Contamination of the main spectra ($s49$ to $s54$) by starlight scattered in the atmosphere should be negligible since the scattered light should not be more than one per cent of the star's flux (Zagury & Goutail 2003; Zagury 2001).

Moving away from the central pixel of a 2-D array, the signal quickly decreases to a spectrum whose shape remains constant (the background, Fig. 1), and is the same on both sides of the central pixel. The background spectrum of Fig. 1 is the average of eight spectra ($s43$ to $s46$, $s57$ to $s60$), taken at a mean distance of $\sim 9''$ from HD 44179.

Backgrounds of the same star, from different observations (top plot of Fig. 4), are similar and differ at most by a slight change of slope (attributed to a difference in atmospheric optical depth). The study of the backgrounds will thus be limited to one per star. The background depends on which star is observed (Fig. 4, bottom plot).

The background slowly decreases with distance from the star so that its precise determination (and an eventual change of shape) is difficult at the edges of the slit. The relationship which exists between the background, the atmosphere, and starlight, is investigated in section 4..

3.2 1-D spectra

A 1-D spectrum, before its absolute calibration, of a point-like object, is determined by adding the spectra of the pixels with larger signal (for the March 2003 observation of HD 44179 (Fig. 1), $s49$ to $s55$ were added).

Top panel of Fig. 5 is a plot of the 1-D spectra, before their final calibration, of all the star observations used in this paper.

1-D spectra of the same object observed at different dates (top panel of Fig. 5) are related to each other by a transformation involving Rayleigh extinction and ozone absorption. On the bottom plot of Fig. 5, second December 2001 and first March 2003 observations of HD 23302 reproduce perfectly well the first December 2001 observation of the star, after multiplication by $e^{-0.015\lambda^{-4}}$ and $e^{-0.0006\lambda^{-4}}e^{-0.6 \cdot 10^{19}\sigma}$ (σ is the ozone cross-section, expressed in cm^2). The corrections to apply to the second December 2001, February 2002, and first March 2003 observations of HD 44179, to match the first December 2001 spectrum are: $e^{0.005\lambda^{-4}}e^{0.5 \cdot 10^{19}\sigma}$, $e^{-0.007\lambda^{-4}}e^{0.5 \cdot 10^{19}\sigma}$, $e^{0.012\lambda^{-4}}$.

3.3 Extended objects

Extended objects are four nebulae and one galaxy. Their size is in the range $15''$ – $25''$, much smaller than the angular size of the slit samples on the sky, but enough to encompass several pixels of the slit.

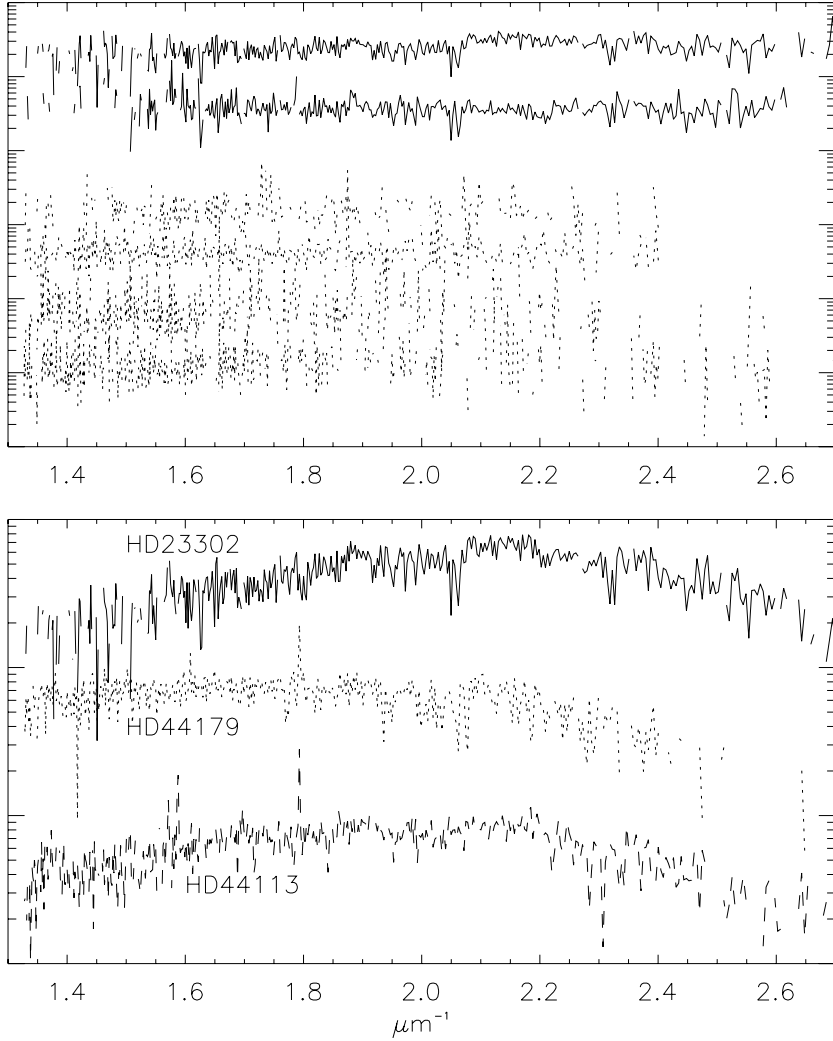


Figure 4. Top: ratios of backgrounds for different observations of HD 44179 (dots, ratios of all observations to the first December one) and HD 23302 (solid lines). The scaling is arbitrary. Bottom: backgrounds (on an arbitrary scale) of HD 44113 (dashes), HD 44179 (dots), and HD 23302 (solid line), observations.

The slit for the Red Rectangle nebula was positioned $\sim 11.5''$ north of HD 44179. The spectra far above the background have the well known bump shape in their near infrared part (Fig. 6, bottom plot). The spectrum with maximum signal is $s59$. Nearest spectra to $s59$ decrease short or long wavelengths first, according to which side of $s48$ the spectrum is, as in the observations of stars (section 3.1). On a larger scale, the decrease of the spectra, from $s59$ to the background, is smooth, and does not show the strong variations found in the observation of stars. The shape of the spectra differs from that of the background, at least in its red part, from $s51$ to $s72$, which represents $\sim 26''$ on the sky, in agreement with the spatial extent of the nebula ($26\text{--}27''$), $11''$

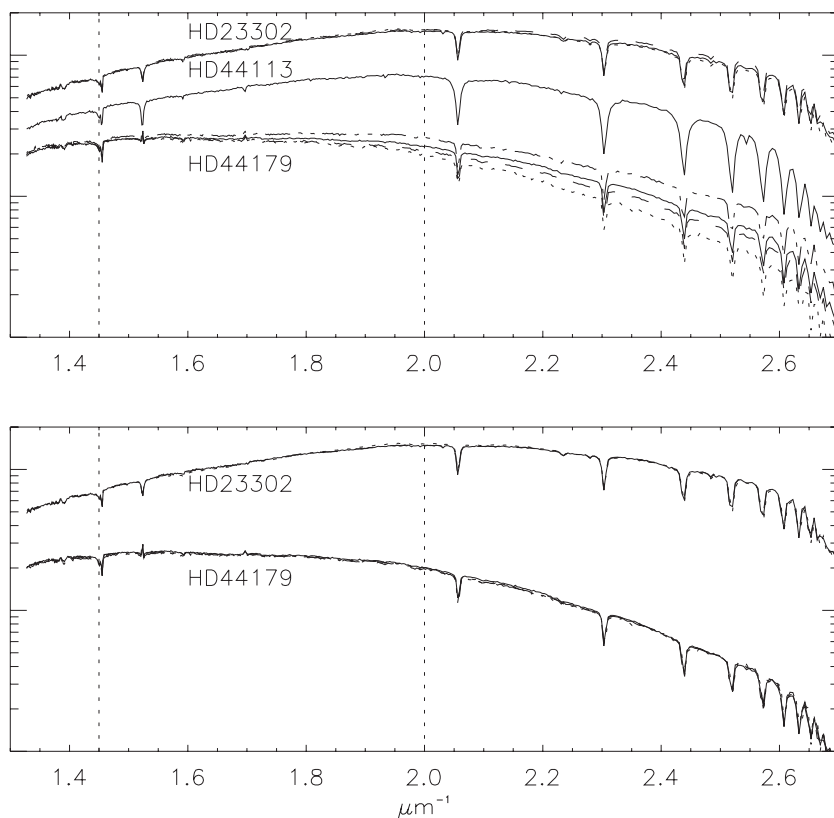


Figure 5. Top: 1-D spectra, before atmospheric correction, of HD 44179, HD 44113, and of HD 23302 observations. Arbitrary scaling. Bottom: 1-D spectra of HD 44179 (resp. HD 23302), from all observations, superimpose well after a correction for atmospheric extinction has been applied.

north of HD 44179. The background, taken at the edges of the slit, is clearly different from the background found in the vicinity of HD 44179.

NGC 2022 was in principle observed 6.5'' south of the star (Karen Kwitter, private communication). The header of the 2-D spectrum fits file, however, which normally indicates the position of the slit, gives an offset of 16'' (this would put the slit on the far-edge, eventually out of the nebula). All spectra from NGC 2022 2-D array are proportional longward of $1.8 \mu\text{m}^{-1}$, with a progressive decrease on each side of the main pixel (pixel 205, Fig. 6). Moving away from s_{205} , a bump-like feature progressively appears upon the spectra, along with the decrease of the spectrum, between $1.5 \mu\text{m}^{-1}$ and $1.8 \mu\text{m}^{-1}$. In none of the spectra, on the possible extent (at most 19'') of the nebula around the direction given by pixel 205, is it possible to see any specific feature, in the continuum, that could be attributed to the nebula. It follows that the spectra, to the precision of these observations, contain no signal from the nebula.

The spectrum of galaxy UGC 11917 (Fig. 7) is representative of many spectra of faint extended galaxies observed with FAST (see the FAST data-base at: <http://tdc-www.harvard.edu/archive/>), which all have a similar cut in intensity around $1.7 \mu\text{m}^{-1}$. The light from UGC 11917 is faint, just above the background. It spreads

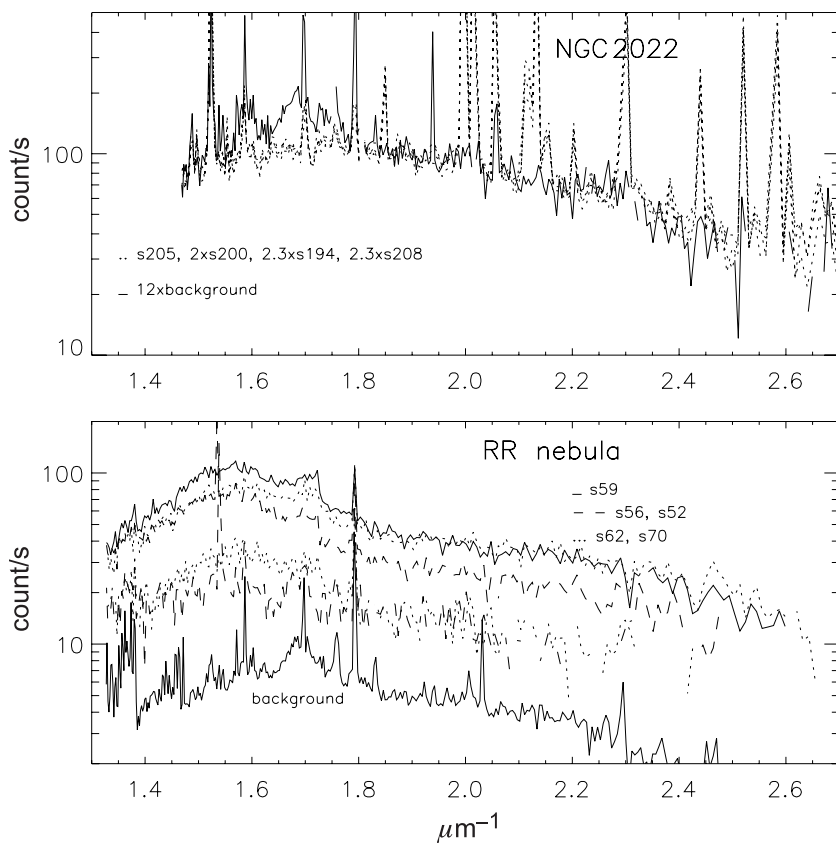


Figure 6. Spectra from the 2-D arrays of NGC 2022 and of the Red Rectangle nebula.

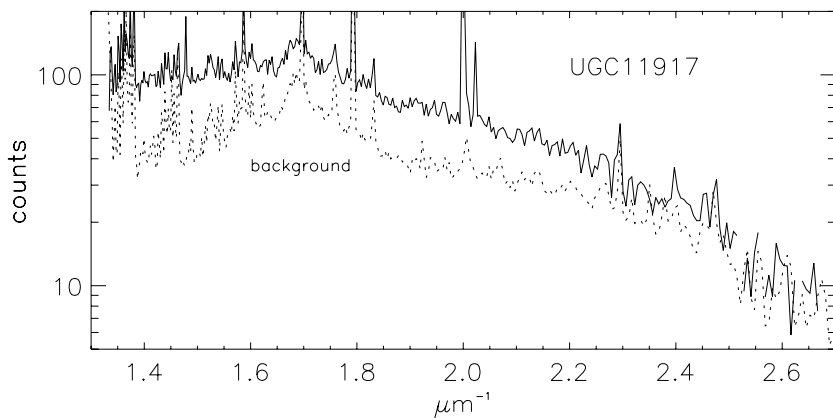


Figure 7. Spectrum s69 of the 2-D array of galaxy UGC 11917, and the background.

$\sim 20''$ (the UGC catalogue gives a minor axis of UGC 11917 of $24''$ in the red) on the sky. There are three identical spectra, *s68*, *s69*, *s70*, on each side of which the signal decreases to the background. Longward of $\sim 2.2 \mu\text{m}^{-1}$ (shortward of $4500 \text{ \AA}$),

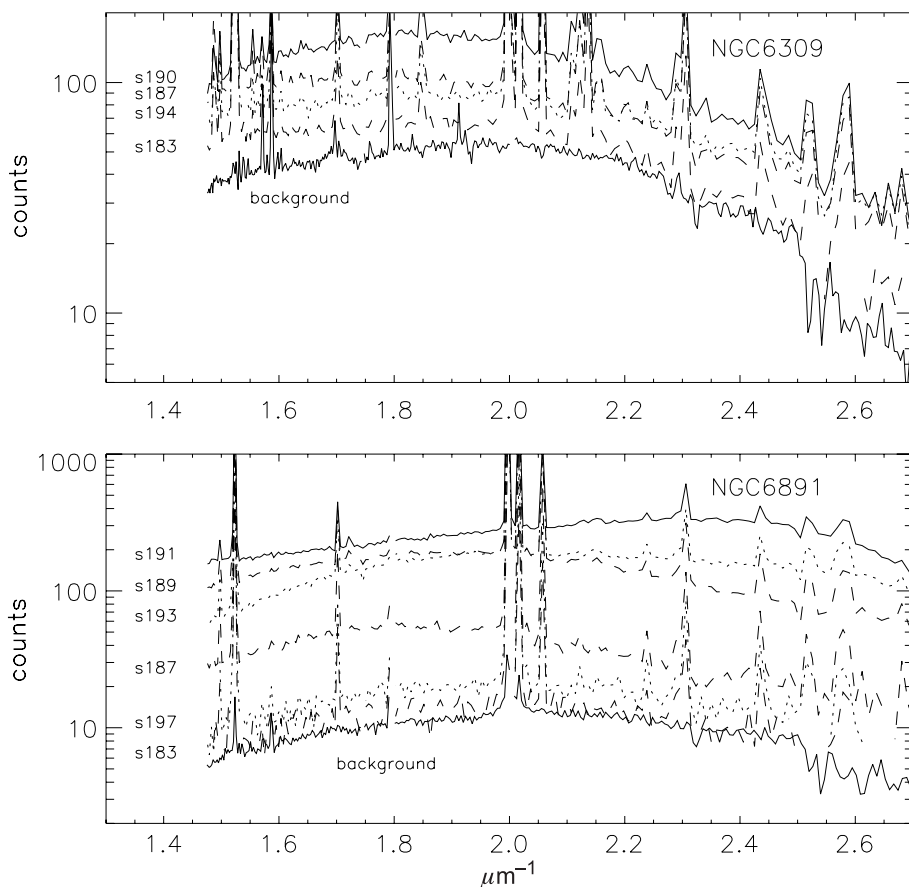


Figure 8. Spectra extracted from the 2-D arrays of NGC 6309 and NGC 6891.

the galaxy spectra meet the background, meaning that, in the near-UV, the galaxy is nearly totally extinguished by the atmosphere.

The slit in NGC 6309 and NGC 6891 observations contains or is close to the central star (Table 3); both 2-D arrays will most certainly include spectra dominated by light from the central star.

The main spectrum for the 2-D array of NGC 6309 is *s*190 (Fig. 8). The shape of the spectrum undergoes strong variations over five pixels (4'') on each side of *s*190 (*s*185 to *s*195), with an increase of the blue slope with decreasing pixel number (as in the star observations). The shapes of *s*184 to *s*182 and of *s*197 to *s*199 differ from the background only in their reddest part ($1/\lambda < 2 \mu\text{m}^{-1}$, Fig. 9). *s*182 and *s*181, *s*200 and *s*201 are nearly identical to the background, which is reached at pixels 180 and 202. The signal is decreased by four between *s*190 and the background.

The main spectrum of NGC 6891 is *s*191 (Fig. 8). The background, ~ 40 times less than *s*191, is reached at pixels 199 and 180. From *s*191 to *s*194, the spectra decrease in the red more quickly than in the blue; *s*195 to *s*199 are proportional to the background. From *s*191 to smaller pixel numbers (pixels 190 and 189), the decrease is

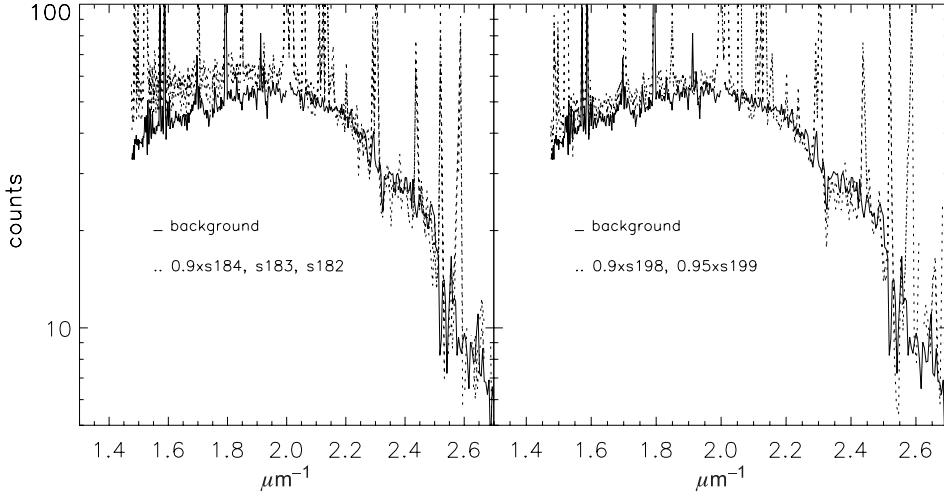


Figure 9. Relation between the background and spectra $s184, s183, s182$ (left plot), $s198, s199$ (right plot), from NGC 6309 observation. An offset of 5 counts was subtracted to all spectra. This offset, with negligible effect in the red, is necessary for the spectra to match the background for $1/\lambda > 2.5 \mu\text{m}^{-1}$. It may correspond to blue light from the nebula since it has to be subtracted on a spatial extent which agrees with the size of NGC 6309.

first more pronounced in the blue. The blue spectra ($1/\lambda > 2 \mu\text{m}^{-1}$) of $s189$ to $s186$ are proportional to the background, the spectra differ in the red. Spectra 184 to 181 are nearly proportional to the background, with a remaining, but small, excess in the red.

In the blue, the pixel range over which the shape of the spectra of NGC 6891 and NGC 6309 differs from that of the background (spectra $s185$ to $s194$ for NGC 6891, $s185$ to $s195$ for NGC 6309) represents ~ 10 pixels ($\sim 8''$), which is half what is expected from the angular extent of the nebulae ($\sim 15''$ or ~ 20 pixels). In the red, NGC 6891 and NGC 6309 differ from the background over $\sim 15''$ (pixels 181 to 198 for NGC 6891, pixels 182 to 199 for NGC 6309), which is about the size of the nebulae. We conclude that nebular light from these two observations is restricted to – at most – a red excess, and a possible small offset in the blue (Fig. 9).

4. Analysis of the background in the observation of stars

Because of the noise level, it is difficult to make a distinction between individual background spectra in the observation of a star. However, adding a sufficient number of these spectra around $10''$ and $65''$ from the direction of HD 23302, a difference will be perceived which proves that the background, if it keeps a constant shape, diminishes with increasing distance from the direction of the star (Fig. 10, left). Adding even more spectra at the edges of the slit for HD 44113, we see (Fig. 10, right) that the shape will also ultimately change.

The straightforward interpretation of the continuum of the background spectra is that it is starlight scattered in the atmosphere, which explains the decrease of the background with angular distance from the star. Following Zagury & Goutail (2003), since the

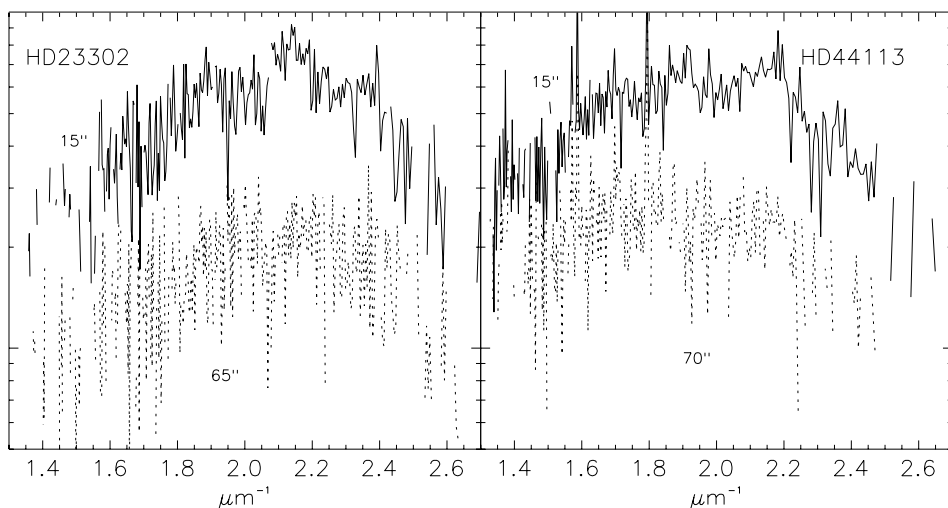


Figure 10. Left is an average of 10 and 40 backgrounds at mean distances 15'' and 65'' from HD 23302. Right is an average of 10 and 20 backgrounds at mean distances 15'' and 70'' from HD 44113. The background for HD 23302 decreases but keeps an identical shape through out the 2-D array. In the backgrounds from HD 44113 observation, one can see the change of shape after the decrease in intensity.

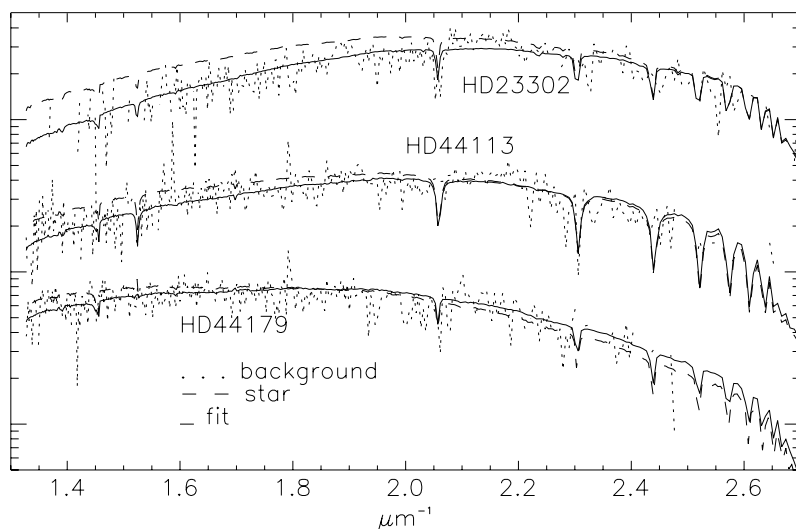


Figure 11. Fit of the star observations backgrounds. Dots: background spectra (first December 2001 observations for HD 44179 and HD 23302, February 2002 observation for HD 44113). Dashes: 1-D spectra of the star (scaled by a constant factor). Solid lines: spectra of the stars multiplied by $e^{-a/\lambda}$ (and a constant factor). a is 0.22, 0.2, 0.1 for HD 44179, HD 44113, HD 23302.

spectra are observed close to the direction of the stars, the scattering is expected to be forward scattering by stratospheric aerosols, with an analytical dependence of the spectra as $1/\lambda e^{-a/\lambda}$. Figure 11 gives the fit found for each background.

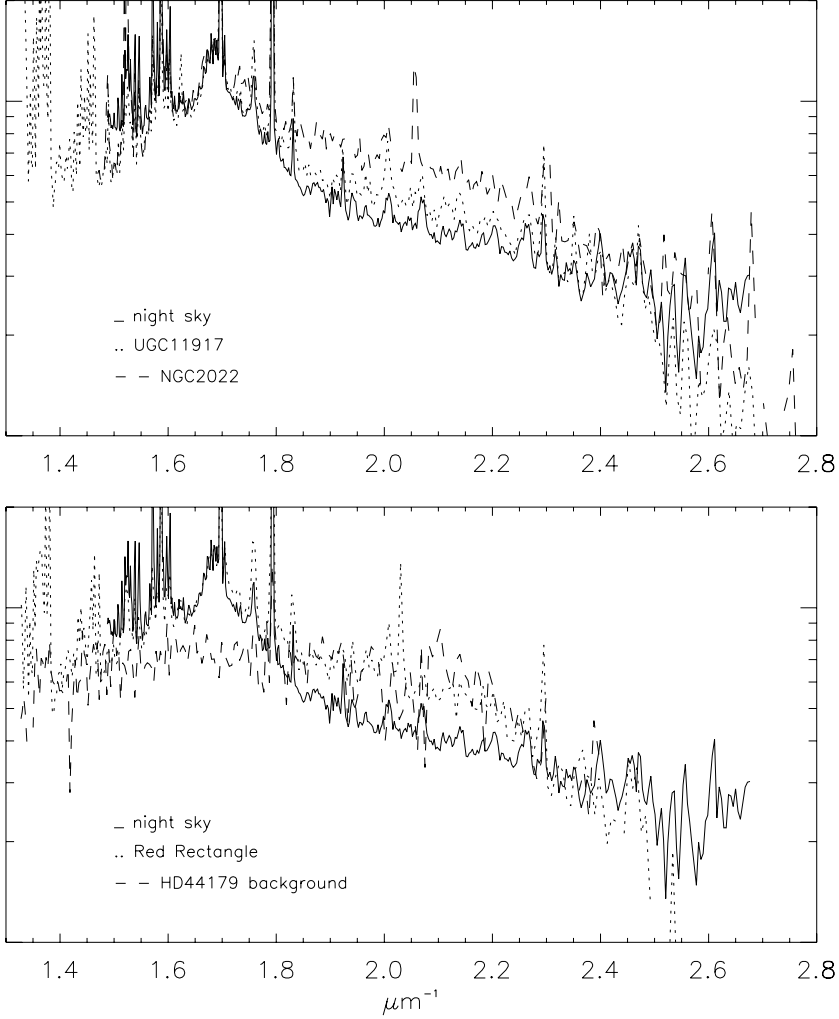


Figure 12. Comparison of extended objects backgrounds to Massey & Foltz (2000) night sky spectrum. Upper panel shows that UGC 11917's background nearly equals the night sky, while there is an excess in the blue for the background of NGC 2022 observation. For the Red Rectangle nebula, this excess corresponds to the background found in the observation of the star and can be attributed to starlight scattered in the atmosphere (lower panel).

5. The background in the observation of extended objects

The backgrounds in the Red Rectangle nebula, in NGC 2022, and in galaxy UGC 11917 are proportional in the red, where they coincide with the night sky spectrum of Massey & Foltz (2000) (Fig. 12).

In the blue, UGC 11917 background nearly equals the night sky spectrum, while NGC 2022 and the Red Rectangle nebula are in excess. In the case of the Red Rectangle, comparison with observations of HD 44179 shows that the blue excess is well fitted by HD 44179 observations background (Fig. 12, lower panel), i.e., it is attributable to light from HD 44179 scattered in the atmosphere.

We deduce that in the observations of UGC 11917, NGC 2022, and the Red Rectangle, the background has reached the night sky spectrum, apart, for NGC 2022 and for the Red Rectangle, for an excess in the blue, due to light from the central star scattered in the atmosphere.

The observation of NGC 2022 is then understood as being mainly light from HD 37882 scattered in the atmosphere for spectrum *s*205, which must correspond to the closest position to HD 37882 on the slit. Moving away from pixel 205, the night sky spectrum progressively appears when the scattered light component diminishes.

The same diminution of the scattered light is observed between the background spectra of the Red Rectangle nebula close to the nebula, and at the edges of the slit. Using the complete set of FAST observations, this will clearly be established in the following article on the Red Rectangle.

The background in NGC 6309 and NGC 6891 observations is not the night sky, and can only be light from the central stars scattered in the atmosphere. Spectra from these 2-D arrays are all (except for the small excess found in the red for some of the spectra) light from the central stars refracted or scattered in the atmosphere.

6. Discussion

The analysis conducted so far shows that the action of atmospheric extinction on starlight can be observed up to more than one arcminute from the direction of a star: direct light from a star, refracted in the atmosphere, and scattered starlight by aerosols, will be felt until their level passes under that of the night sky.

Over a radius of a few arcseconds (3–8'' for the observations presented here) around the direction of the star, direct light from the star, which has followed different paths through the atmosphere, dominates the spectrum. In this angular region, the orientation of the slit in the azimuth-altitude plane probably determines the variations of the shape of the spectrum.

After this central angular region, the spectrum has decreased by a factor of ~ 100 ; starlight forward-scattered by aerosols in the atmosphere contributes most. The shape of the spectrum, which is proportional to the spectrum of the star times $1/\lambda$, remains constant. The distribution of the scattered light is spherically symmetric (it is the same on both sides of the slit when moving away from the star). Its intensity decreases with distance from the star. Starlight scattered by aerosols will ultimately fade under the background night sky. The ratio of scattered starlight (which depends on the star magnitude) to night sky (independent of the star), as a function of distance from the star, should be fixed by the magnitude of the star.

If the star illuminates a nebula, lights scattered by the nebula and by the atmosphere will compete in the observed spectrum, and, at small angles, with direct light from the star. Contrary to the ratio of starlight scattered in the atmosphere to night sky, the ratio of starlight scattered in the atmosphere to nebular light should not depend on the magnitude of the illuminating source. It will depend on the degree of atmospheric extinction (thus on wavelength, on airmass, and on atmospheric conditions), since, if atmospheric extinction is low, light from the nebula will more easily reach the observer, while starlight scattered in the atmosphere will be diminished.

It is inferred that nebular light will more easily be perceived in the red, where atmospheric extinction is minimized, rather than towards the UV, where Rayleigh atmospheric extinction ($\propto 1/\lambda^4$) is important. On the other hand, starlight scattered in the

atmosphere will first appear in the blue, also because atmospheric extinction (and therefore atmospheric scattering by aerosols) is enhanced when moving to UV wavelengths. In between these two wavelengths' domains one finds the ozone absorption band.

The continua from the 2-D arrays of the observations of nebulae examined in section 3.3 provide a direct application of these conclusions. Nebular light is obvious for the Red Rectangle in the red, though in the blue the present analysis does not permit to separate the spectrum of the nebula from that of light from HD 44179 scattered in the atmosphere, and from the night sky. For NGC 6309 and NGC 6891 spectra, the part due to the nebula is small, restricted to an excess in the reddest part of the spectrum, eventually to an offset in the blue. In these observations, the proximity of the central star, and its extinction by the atmosphere, clearly determine the shape of the continuum, in and out of the nebula. No nebular light was detected for NGC 2022: the decrease of the spectra on each side of pixel 205 illustrates the passage from starlight scattered in the atmosphere to night sky spectrum.

Close to the direction of a star forward scattering by large grains (aerosols for the atmosphere) should prevail both for the light scattered by a nebula and for the light scattered by the atmosphere, with an analytical dependence $\propto 1/\lambda e^{-a/\lambda}$, as it was found for the backgrounds in the observation of stars. The angle of scattering for starlight scattered in the atmosphere will always be extremely small. This may not be the case for starlight scattered by nebulae, since the angle of scattering can be large if the nebula is close to its illuminating star. Under these conditions, Rayleigh scattering by the gas could prevail in some nebulae, with a λ^{-4} (rather than λ^{-1}) dependence on λ .

7. Conclusion

This paper has considered different data-sets issued from long-slit, ground-based, spectroscopy at visible wavelengths, of stars, nebulae, and of one galaxy (although representative of faint extended galaxies observed with FAST). The objects were separated into two classes: point-like objects, as stars, and extended objects. The paper has surveyed different aspects of the influence on these observations atmospheric extinction has, and it has lingered over the study of the background (i.e., the signal away from the object) spectra in these observations.

The study of the 2-D spectra of stars has shown that the direct light from a point-source covers a few arcseconds of the slit. By a clear sky, stars' spectra are little affected by ozone absorption, while Rayleigh extinction causes the important bending observed at the shortest wavelengths ($1/\lambda > 2 \mu\text{m}^{-1}$; $\lambda < 5000 \text{ \AA}$). At larger distances, the spectrum is that of starlight scattered by aerosols in the atmosphere, with a wavelength dependence as $e^{-a/\lambda}/\lambda$ (and proportional to the spectrum of the star). The spectrum of the scattered light will decrease slowly, and keep a steady shape, until it becomes negligible compared to the night sky spectrum.

In the spectra of extended objects, several effects of different nature are superposed. They are therefore more difficult to interpret.

The 2-D spectrum of an extended object results from:

- Light from the region of the extended object itself, sampled by the pixel under consideration, which will more easily be perceived in the red (towards the near-infrared).

- Direct light from the central source (if any) refracted in the atmosphere, which can be efficient up to a few arcseconds from the source.
- Light from the central source scattered in the atmosphere, which can be present over more than one arcminute, mainly in the short wavelength range.
- The night sky, which contributes to the spectrum when the preceding sources of light are faint enough.

It may also be (as for point sources) that light from the nearest regions, refracted by the atmosphere, contributes to what is received by a given pixel. However, since refracted light quickly decreases with distance, this should have a minor effect.

For nebulae, we noted that the relative strength of light from the nebula and starlight scattered in the atmosphere does not depend on the magnitude of the illuminating star (but depends on the degree of atmospheric extinction).

Spectra above the background, in the observations of NGC 6309 and of NGC 6891, were found to be mainly direct light from the central sources, and light from these sources scattered in the atmosphere; the continuum of nebular light in these observations is probably not more than a small excess in the red and an offset in the blue.

The 2-D array of NGC 2022 shows no nebular light and was proved to be light from HD 37882 scattered in the atmosphere for the pixels with the highest signal. This observation illustrates the passage from scattered light in the atmosphere to night sky spectrum when moving away from the central star.

It is only in the observation of the Red Rectangle nebula that nebular light clearly appears in the red. However, the contribution of light from HD 44179 scattered in the atmosphere was not separated in this observation. Additional observations will help, in a companion paper, to precise the relative contribution of each.

It can already be seen that the determination of nebular spectra from ground-based observations can not be reduced to the traditional subtraction of a background followed by calibration by a standard star. In particular, the continuum of ground-based spectra of nebulae presented in the Kwitter–Henry database is not reliable; this may not necessarily affect emission lines (only the level of the continuum is changed), studied by K. Kwitter and collaborators, but should not be neglected.

The data in hand have not permitted to deepen some problems in the data reduction process and on the understanding of the background. The main spectra of the 2-D arrays in star observations are related in a way (see section 3.1, and figures 1 and 3) which would merit attention and a better understanding. The study of the background in long-slit spectroscopy initiated here can be improved by a more precise evaluation of its dependence with stars' angular distance and magnitude.

Acknowledgements

I thank R. Cutri and J. Huchra for their help in obtaining the FAST data. Most of all, I am indebted to Lucas Macri who repeatedly observed the targets used in the paper. Susan Tokarz has kindly reduced the data, and answered many questions on the procedures she uses, as well as on the FAST instrument. I also thank K. Kwitter for agreeing to the release of some of her Kitt Peak observations.

References

- Acker, A. *et al.* 1992, *The Strasbourg/ESO catalogue of Galactic Planetary Nebulae* (Garching bei München: ESO).
- Fabricant, D. *et al.* 1998, *PASP*, **110**, 79.
- Kwitter, K. B., Henry, R. B. C. 2001, *ApJ*, **562**, 804.
- Kwitter, K. B., Henry, R. B. C. 2002, *Bulletin of the AAS*, **34**, 1251.
- Kwitter, K. B., Henry, R. B. C., Milingo, J. B. 2003, *PASP*, **115**, 80.
- Massey, P., Foltz, C. B. 2000, *PASP*, **112**, 566.
- Tokarz, S. P., Roll, J. 1997, *A.S.P. Conference Series*, Gareth Hunt and H. E. Payne (eds), **125**, 140.
- Zagury, F. 2001, *NewA*, **6/7**, 415.
- Zagury, F., Goutail, F. 2003, *NewA*, **8**, 537.
- Zagury, F., Fujii, M. 2003, *NewA*, **8**, 549.

An Analysis of Spectra in the Red Rectangle Nebula

Frédéric Zagury

Institut Louis de Broglie, 23 rue Marsoulan, 75012 Paris, France.

e-mail: fzagury@wanadoo.fr

Received 2004 September 27; accepted 2005 August 24

Abstract. This paper presents an analysis of a series of spectra in the Red Rectangle nebula. Only the reddest part of the spectra can safely be attributed to light from the nebula, and indicates Rayleigh scattering by the gas, in conformity with the large angles of scattering involved and the proximity of the star. In the blue, light from HD 44179, refracted or scattered in the atmosphere, dominates the spectra. This paper questions the reliability of ground-based broad-band spectra of extended objects in the blue.

Key words. Atmospheric effects—planetary nebulae: general—planetary nebulae: individual (HD 44179).

1. Introduction

This paper will pursue the investigations initiated in a preceding article (Zagury 2005) on the effects of atmospheric extinction on starlight.

Long slit spectroscopy of a sample of objects (stars, nebulae, galaxies) was used to determine the nature of the light received by a telescope at close angular distances from a star, before the spectrum reaches the night sky spectrum. Direct light from the star, refracted by the atmosphere, will be detected over the first few arcseconds. Over the next few tens of arcseconds, forward scattering of starlight by aerosols will dominate the spectrum, until it fades out under the night sky spectrum.

Light from the central star, refracted or scattered in the atmosphere, cannot be neglected when dealing with spectra of a nebula, since it will compete with the light scattered by the nebula. The continua of the spectra of NGC 6309, NGC 6891, and NGC 2022, for instance, as they are given in the Kwitter–Henry database (<http://oit.williams.edu/nebulae/browse.cfm>), more likely reflect the spectrum of the central star, than the true nebular spectra (Zagury 2005).

If the preceding paper emphasized the importance of atmospheric extinction for our understanding of ground-based observations of nebulae, it was not quantitative enough to separate, in the observed spectrum of a nebula, atmospheric effects from true nebular spectrum. This problem will guide the present paper, which will focus on spectra of the Red Rectangle nebula, observed at different positions in the nebula, and at different dates.

2. Observations

The observations of the Red Rectangle used in this article come from three observing runs at the Fred L. Whipple Observatory, in December 2001, February 2002, and March

Table 1. Observations.

Position	U.T. ⁽¹⁾	Δt ⁽²⁾	$\Delta \delta$ ⁽³⁾	A.M. ⁽⁴⁾	Alt. ⁽⁵⁾
December 2001 (2001-12-22)					
HD 44179-1	08:44:21	15		1.41	44.92
<i>n1</i>	08:45:08	60	5.4	1.41	44.80
<i>n2</i>	08:46:53	120	9.8	1.42	44.57
<i>n3</i>	08:57:00	240	14.3	1.45	43.42
<i>n4</i>	09:03:31	120	11.5	1.46	42.96
<i>n5</i>	09:07:44	120	7	1.48	42.52
HD 44179-2	09:10:04	15		1.48	42.44
February 2002 (2002-02-09)					
HD 44113	5:05:06	15		1.38	46.28
HD 44179	5:07:24	5		1.37	46.59
<i>n1</i>	5:14:08	60	5.3	1.38	46.15
March 2003 (2003-03-26)					
HD 44179-1	3:10:46	5		1.51	41.40
<i>n1</i>	3:11:46	60	5.4	1.51	41.18
<i>n2</i>	3:13:23	120	9.8	1.52	40.86
<i>n3</i>	3:15:59	240	14.3	1.54	40.30
<i>so1</i>	3:20:33	120	-5.3	1.55	39.97
<i>n1b</i>	3:23:11	240	6.7	1.57	39.36
<i>so2</i>	3:28:09	240	-8.3	1.59	38.71
HD 44179-2	3:33:27	5		1.60	38.52

⁽¹⁾ UT HH:MM:SS at the start of exposure.

⁽²⁾ Duration of exposure (seconds).

⁽³⁾ Elevation offset from HD 44179.

⁽⁴⁾ Air mass.

⁽⁵⁾ Altitude (°) of the object at the time of observation.

2003. Spectra were acquired with the FAST spectrograph, mounted on the Tillinghast 1.5 meter telescope. Technical information on FAST will be found in Fabricant *et al.* (1998).

FAST is a long slit spectrograph, which samples a region 3' long and 3'' wide on the sky and covers (in the configuration used for the observations) the [3660 Å, 7530 Å] wavelength range, with a spatial resolution of 0.61''/pixel. The slit is oriented E–W. Different positions of the slit across the nebula are thus characterized by their declination offset from HD 44179.

The 2-D spectra presented here were binned by two (December 2001, March 2003) or by four (February 2002), along their spatial dimension, during the data reduction process (Tokarz & Roll 1997), yielding a final angular resolution of 1.2'' or 2.4'' per pixel. 1-D spectra of stars are extracted from the 2-D arrays by adding the spectra far above the background. The wavelength resolution is 1.4 Å/pix. The seeing during the three observing runs was good, 1–2''.

The main information on the data is grouped in Table 1. Figure 1 shows the projected positions of the slit on the nebula, for each run, and specifies the position on the slit of the pixel with maximum signal.

An observation will be designated by its position (*n1*, *n2*, ...), as it is given in Table 1, and the date of observation. The spectrum at pixel 'x', extracted from the 2-D array of an observation, will be noted 'sx'. 'σ' stands for the ozone cross-section,

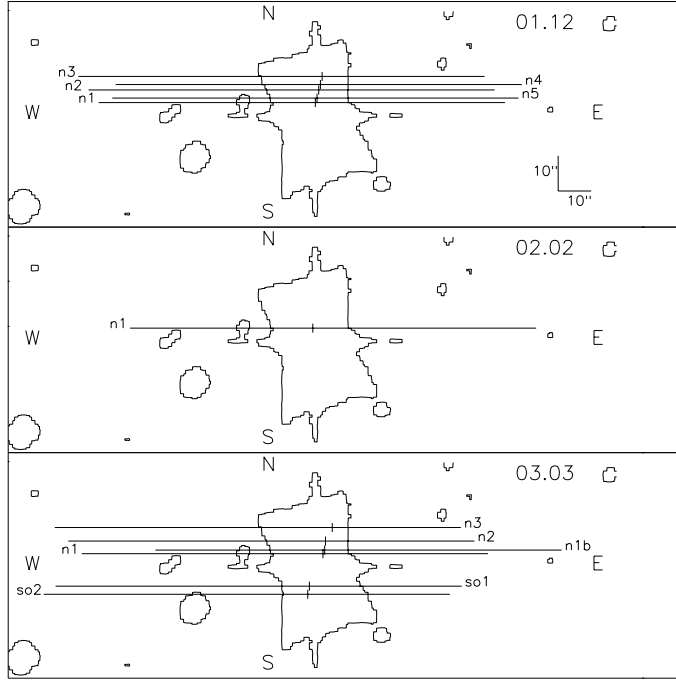


Figure 1. Summary of the observations of the Red Rectangle nebula. For each run (one plot per date), the successive positions of the slit (horizontal lines) are projected on a contour plot of the DSS2 red survey image centered on HD 44179 ($\alpha = 06^h 19^m 37^s$, $\delta = -10^\circ 38' 15''$, J2000). The field is $3.6' \times 1.8'$ large. The slit is oriented E–W, pixel numbers increase with right ascension. On each line, the approximate position of the pixel (of the 2-D array) with maximum signal is marked by a vertical line.

and is expressed in cm^2 . Vertical dotted lines on the plots give the limits of the ozone absorption region.

3. December 2001 observations

3.1 HD 44179

Spectra extracted from the 2-D arrays of HD 44179's observations, at the beginning and at the end of December 2001 run, are plotted Fig. 2.

For each 2-D array a main spectrum is identified ($s61$ for the first observation, $s64$ for the second), which fixes the direction of HD 44179. On each side of the main pixel the intensity decreases steeply, to less than ~ 1 count/s, over ~ 7 pixels or $\sim 8''$. The shape of the spectra with level above 1 count/s evolves in a regular way, with a systematic increase of the average slope of the spectra with increasing pixel number (see Zagury 2005). Some spectra, $s65$ and $s64$ in the first observation, $s60$, $s59$, in the second one, have a slight bump in the red (the red bump is more pronounced in February 2002 observation, section 4.) which is due to the nebula. $8''$ away from the direction of HD 44179, the spectra have reached a shape (mentioned as the background on the plots) which is common to the following spectra and to both sides of HD 44179.

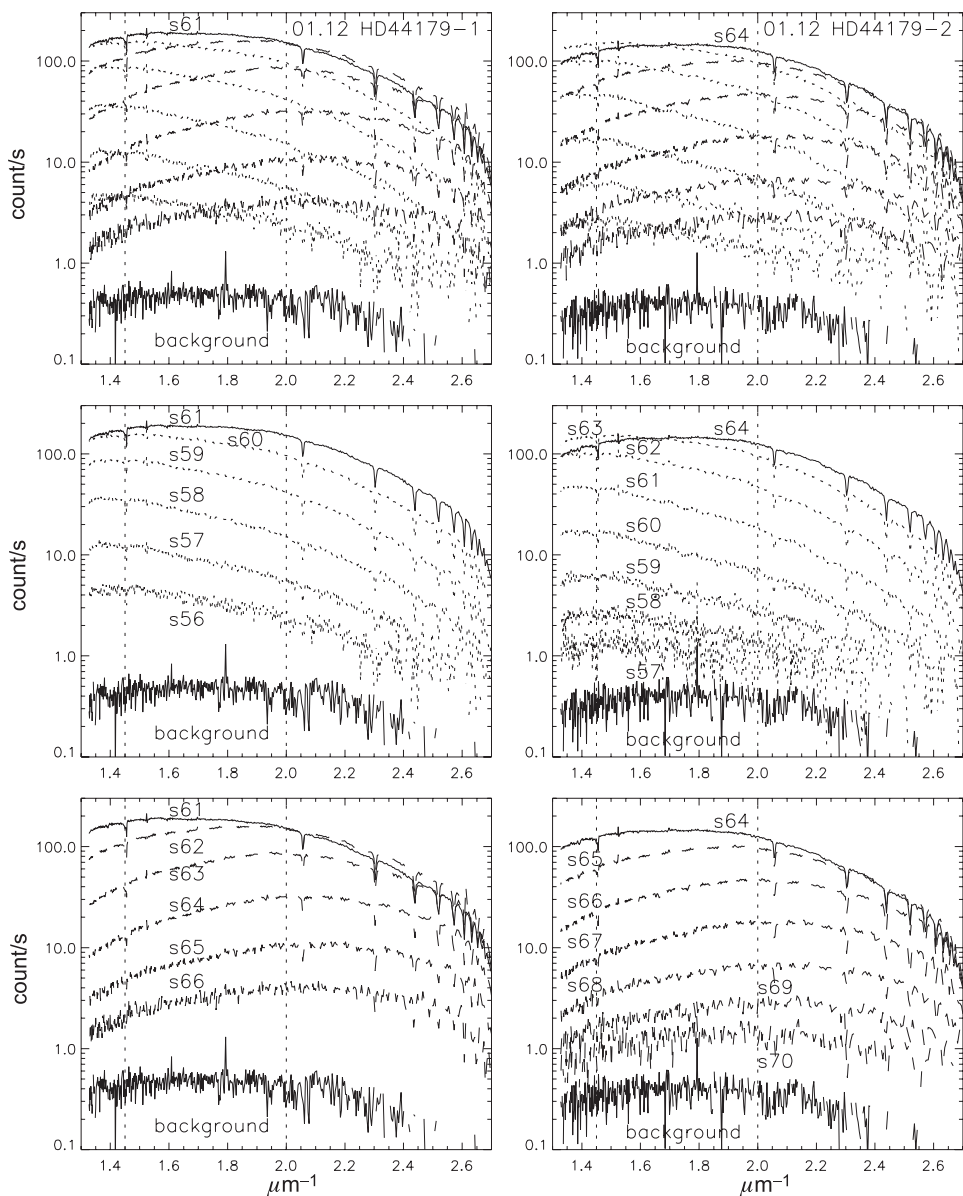


Figure 2. 2-D arrays for the December 2001 first (left column) and second (right column) observations of HD 44179.

The noise level is ~ 0.2 count/s, less but not far from the level of the background. The background spectra of Fig. 2 are the average of 8 spectra at a mean distance of $11''$ from HD 44179. The backgrounds, identical in the two observations, are interpreted (Zagury 2005) as light from HD 44179 scattered by aerosols in the atmosphere, with a spectrum proportional to the spectrum of HD 44179 and to $1/\lambda$ (Fig. 3). Moving away from HD 44179, the background progressively diminishes in intensity, with a

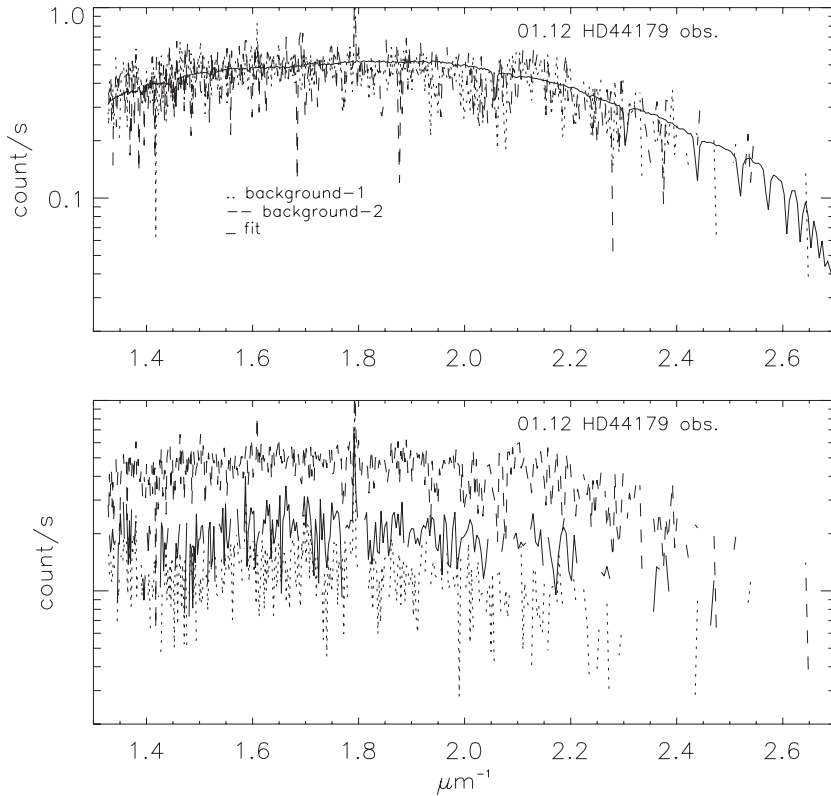


Figure 3. Upper panel: Backgrounds (average of eight spectra of the 2-D arrays, at a mean distance of $11''$ from HD 44179) for the two observations of HD 44179 have an identical shape, proportional to the spectrum of HD 44179 times $1/\lambda$. Lower panel: Decrease of the first observation background with distance from HD 44179. The backgrounds are the average of 8, 22, 40 spectra, taken at mean distances of $11''$, $30''$, and $60''$, from HD 44179.

constant shape, which will ultimately also be modified, because of the attenuation of the scattering and the merging of the night sky spectrum (Fig. 3, bottom plot).

Between the beginning and the end of the run the airmass has increased from 1.41 to 1.48. The result is a decrease of the 1-D spectrum of HD 44179 by a factor of 1.15, a slightly more pronounced depression in the ozone absorption region, and an increase of the bending in the blue (a decrease of the blue slope, bottom spectra of Fig. 4). These differences between the spectra can be corrected analytically (top spectra of Fig. 4).

3.2 Nebula

The 2-D arrays of the December 2001 nebular spectra are presented in Figs. 5 (right column) and 6. Spectra from positions $n4$ and $n5$, which are similar to $n3$'s (see below), were not represented. The main spectrum for each observation is in plain line. Its position on the slit, and in the nebula, is marked on Fig. 1.

The red bump, characteristic of the Red Rectangle, is present in all 2-D arrays, with a spatial extent that matches the width of the nebula at the corresponding declination (derived from the red Digital Sky Survey (DSS2) image).

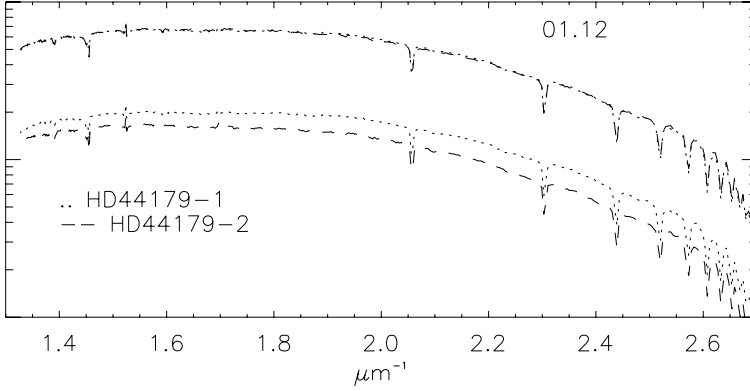


Figure 4. Comparison of the two December 2001 1-D spectra of HD 44179. Bottom spectra: Between the beginning (dots) and the end (dashes) of the run, the 1-D spectrum of HD 44179 has decreased in intensity. The slope of the spectrum is slightly decreased in the blue. Top spectra: The difference of atmospheric extinction between the two observations is corrected by multiplying the second observation 1-D spectrum by $1.15e^{0.005/\lambda^4} e^{8 \cdot 10^{18} \sigma}$.

Due to the long exposure times low level background spectra can be determined with better accuracy than for the 2-D arrays of HD 44179. Two background spectra are represented in solid line on each plot, one immediately outside the nebula, one close to the extremities of the slit. Backgrounds of all nebular observations, taken at the extremities of the slit, are identical (in shape and intensity, Fig. 7, left). Their shape matches the night sky spectrum of Massey & Foltz (2000), except in the blue where there is an excess of signal for the backgrounds (Fig. 7, middle). This excess is attributed to remaining light from HD 44179 scattered in the atmosphere, and is well fitted by the 1-D spectrum of HD 44179 (first observation of the run) times $1/\lambda$ (and a constant factor, middle plot of Fig. 7). Subtraction of the scattered light component will reproduce the Kitt Peak night sky spectrum (right plot of Fig. 7). Backgrounds taken close to the nebula are equal to the far-away backgrounds with an additional component of light from HD 44179 scattered in the atmosphere (Fig. 8).

The first position observed in the nebula, $\sim 5''$ north from HD 44179, is close enough to HD 44179 to consider a possible contamination of the spectra by direct light from HD 44179 (as for NGC 6309 and NGC 6891 in Zagury (2005)). In the blue, the main spectrum (s56, right plot of Fig. 5) of the 2-D array recalls, by its shape as well as by its intensity, spectra of the 2-D arrays of HD 44179 taken at approximately equal distances from the star, s64 and s65 of HD 44179 first observation, s66 and s67 of the second observation (Fig. 9). It is (in the blue) exactly proportional to s65 of the first observation, and nearly proportional to s67 of the second one (a small correction for gas extinction is necessary to superimpose the spectra in the blue); the spectrum of the nebula is superimposed, in the red, on the spectrum of HD 44179.

Outside the nebula, still at position *n* 1, the shape of the background spectrum evolves from a predominance of light from HD 44179 scattered in the atmosphere (top background spectra of Fig. 5, right column plots, with a shape identical to the background found for the observation of HD 44179) to a predominance of the night sky spectrum (bottom background spectra), at the edges of the slit.

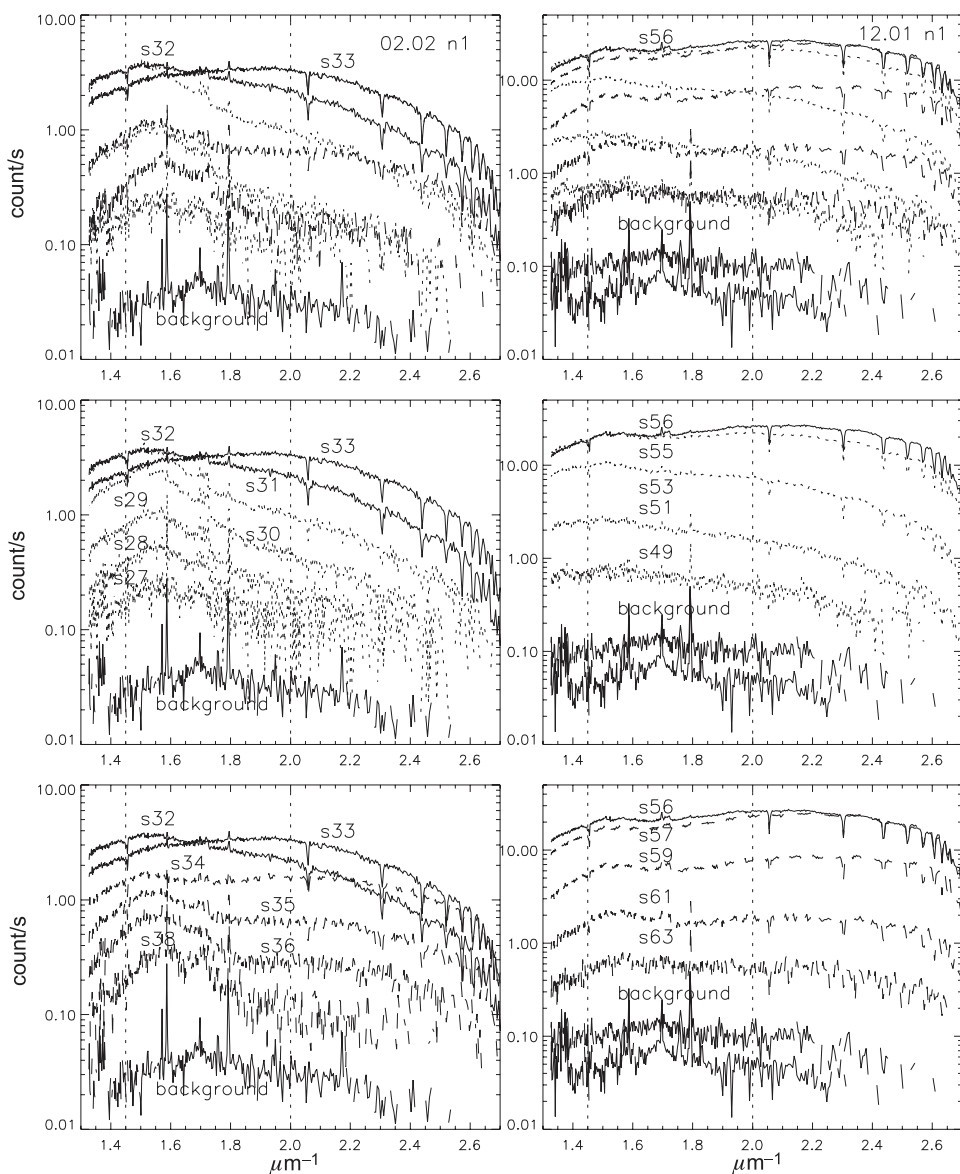


Figure 5. December 2001 (right) and February 2002 (left) observations of position *n1*. Two backgrounds, close to the nebula and at the edges of the slit, are plotted for the December 2001 observation, which show the transition from HD 44179's light scattered in the atmosphere to night sky spectrum.

The other observations of the nebula are farther away from HD 44179; the proportion of direct light from HD 44179 should be highly diminished, which explains the important fall in intensity of the spectra (compare Figs. 6, 9, and 10).

Main spectra at positions 3, 4 and 5 have rigorously identical shapes; they differ by a constant factor close to 1 (dotted spectra of Fig. 10). The blue part of *n3*, *n4*, *n5*, 2-D arrays keeps a constant shape in and out of the nebula (Fig. 11), with only a diminution

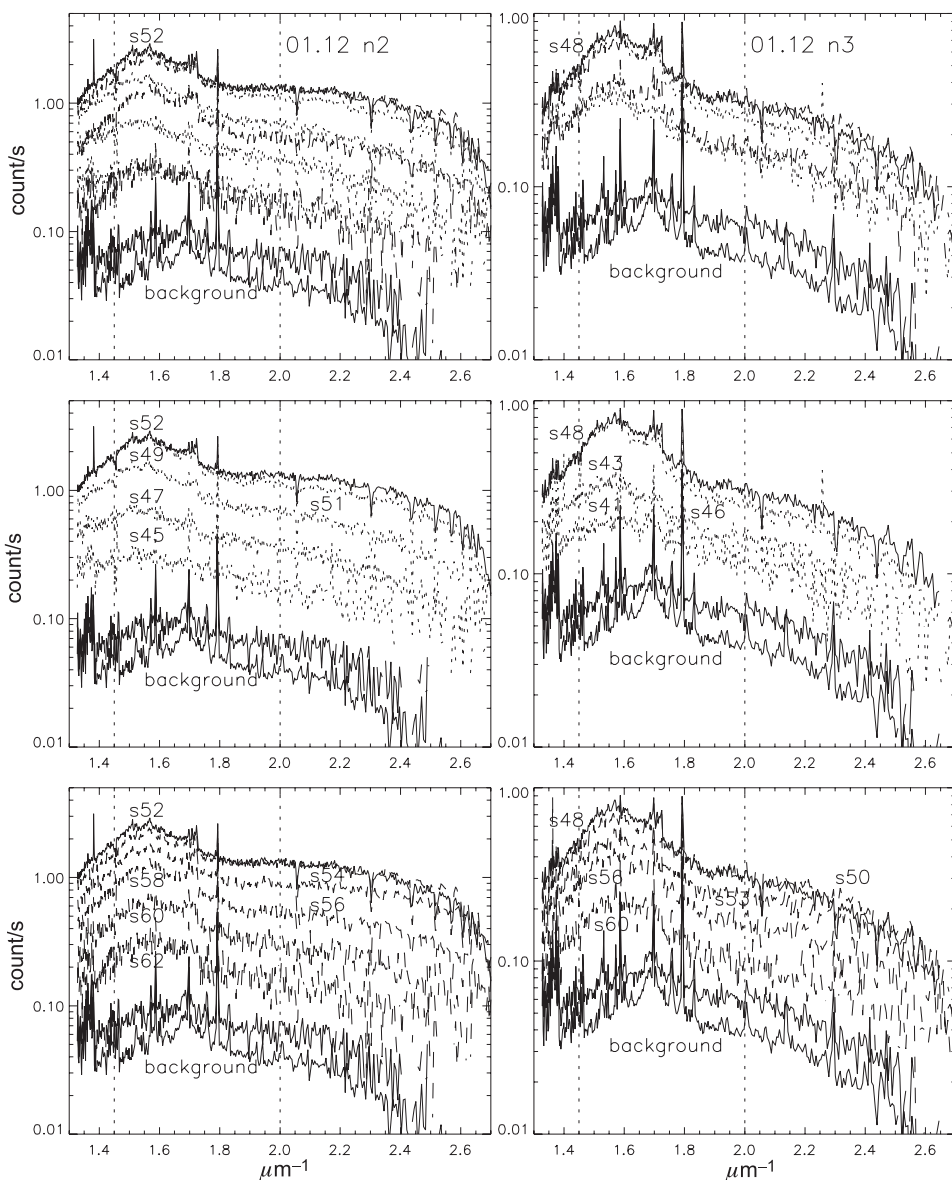


Figure 6. December 2001 observations of the nebula at positions *n2* (left) and *n3* (right).

of the magnitude when moving away from it. The spectra are also observed at about the same distance ($8''$ – $15''$) from HD 44179 where the backgrounds for the HD 44179 observations were established (section 3.1). In the blue they are similar to these backgrounds, both in shape and in magnitude (a few 0.1 count/s). The backgrounds from the observations of HD 44179 would superimpose well, with a scaling factor close to 1, on the blue part of the spectra of Fig. 11. On the figure, the backgrounds are fitted, in the blue, by the spectrum of scattered light (by atmospheric aerosols) from HD 44179 (proportional to the product of the spectrum of HD 44179 by $1/\lambda$).

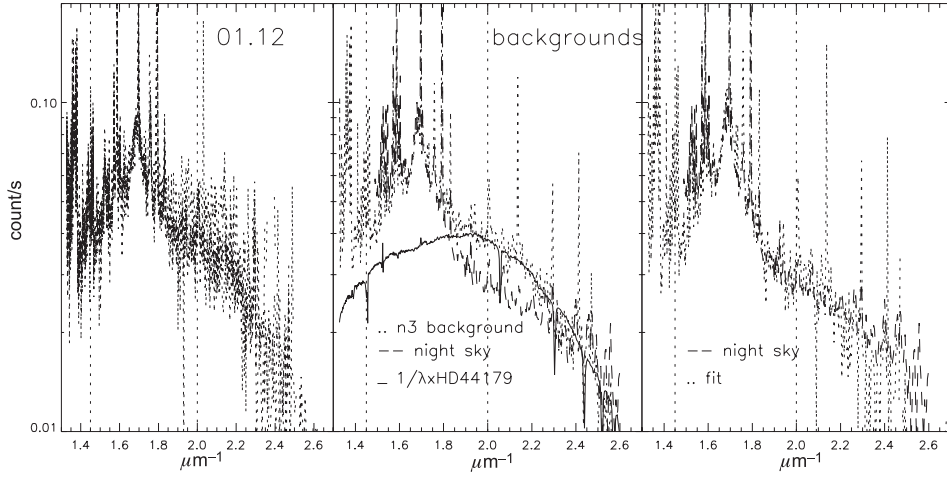


Figure 7. Left: Backgrounds, taken at the extremities of the slit, from December 2001 observations of the nebula ($n1$ to $n5$), are all identical. Middle: These backgrounds correspond to the night sky spectrum (dashes, from Massey & Foltz 2000), superimposed on an additional component of light from HD 44179 scattered in the atmosphere (proportional to the spectrum of HD 44179 multiplied by $1/\lambda$), which mainly affects the blue part of the spectrum. Right: The Kitt Peak night sky is reproduced by the difference between the background of the nebular observation and a component of light from HD 44179 scattered in the atmosphere.

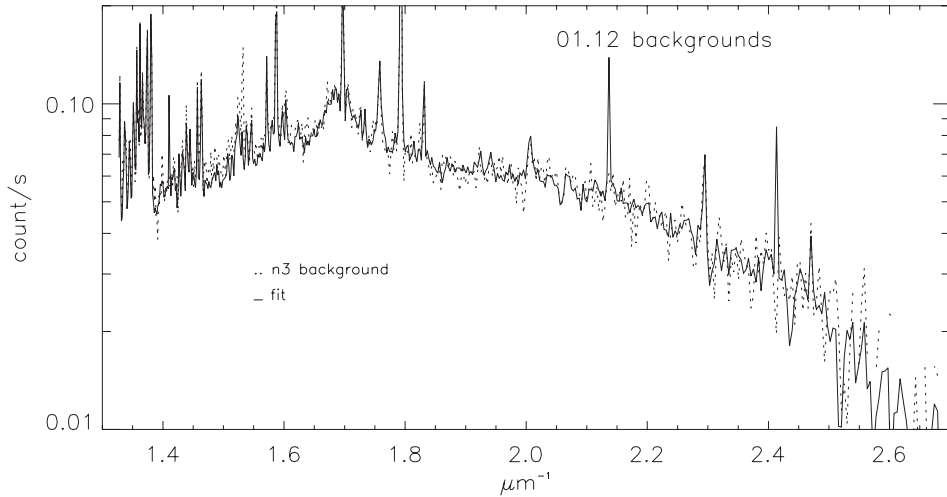


Figure 8. The backgrounds close to the nebula are deduced from the backgrounds away from it by an additional component of light from HD 44179 scattered in the atmosphere. Here, the background, close to the nebula, for the December 2001 observation at position $n3$, is fitted by the background at the extremities of the slit, added to the 1-D spectrum of HD 44179 multiplied by $1.2 \cdot 10^{-6}/\lambda$.

Spectra from $n3$, $n4$, $n5$, 2-D arrays, in the nebula, are thus made of two components: nebular light in the red, stacked on scattered light (from HD 44179) in the atmosphere, which dominates in the blue. The red rise is constant for the closest positions to

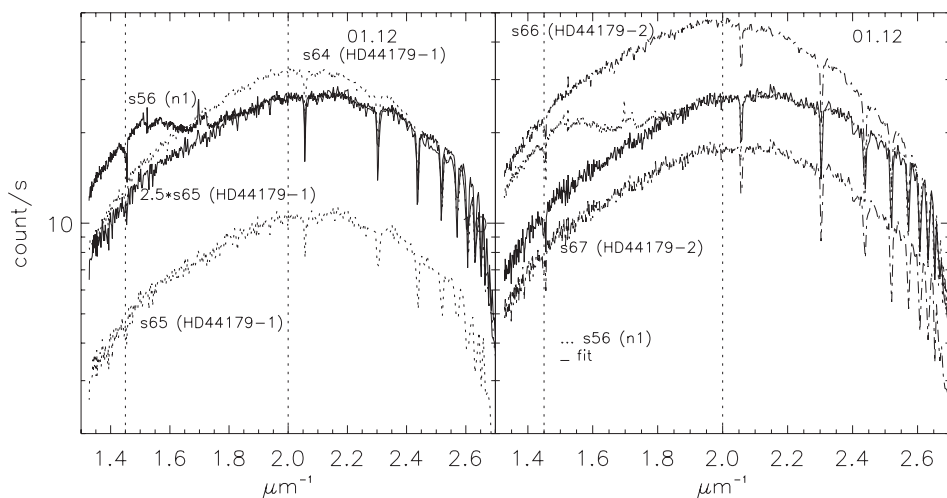


Figure 9. Left: Main spectrum ($s56$) of $n1$ 2-D array is in-between $s64$ and $s65$ of the first observation of HD 44179, and is in the blue well reproduced by $2.5 \times s65$. Right: $s56$ of $n1$ is compared to spectra of the second observation of HD 44179. It is in-between $s66$ and $s67$, and nearly proportional to $s67$ (the fit is $1.3e^{0.0075/\lambda^4} s67$). The spectrum of the nebula appears in the red only, superposed on the spectrum of HD 44179.

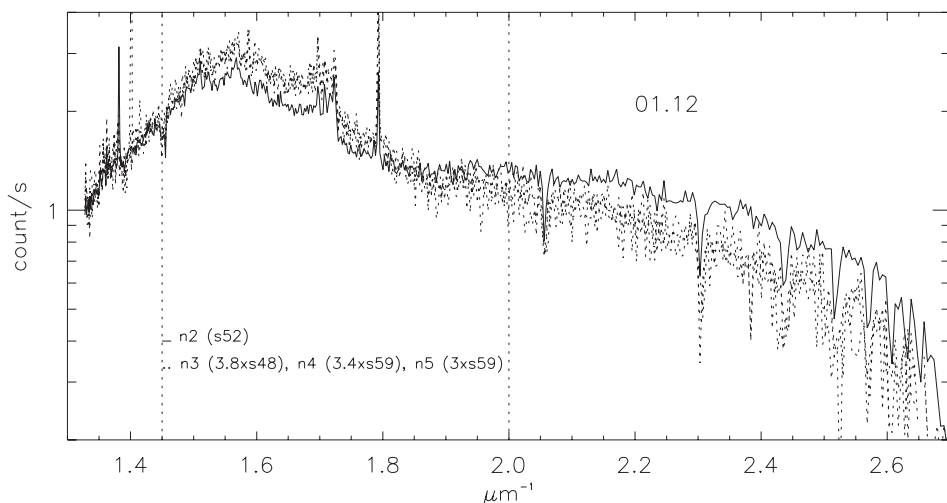


Figure 10. Comparison of the main spectra of December 2001 observations of the nebula. Positions 3, 4 and 5, in dots, differ by a constant factor only.

the main pixel, but its slope will diminish towards the edges of the nebula, because of the increasing proportion, in the spectra, of scattered light in the atmosphere.

The main spectrum ($s52$) of the 2-D array at position $n2$ is of a higher level. The red rise is as for the main spectra at positions 3, 4, and 5, but the spectra differ in the blue (Fig. 10). The blue part of $s52$ is in-between, and proportional to spectra $s67$ and $s68$, from the first observation of HD 44179 (Fig. 12), observed at about the same angular

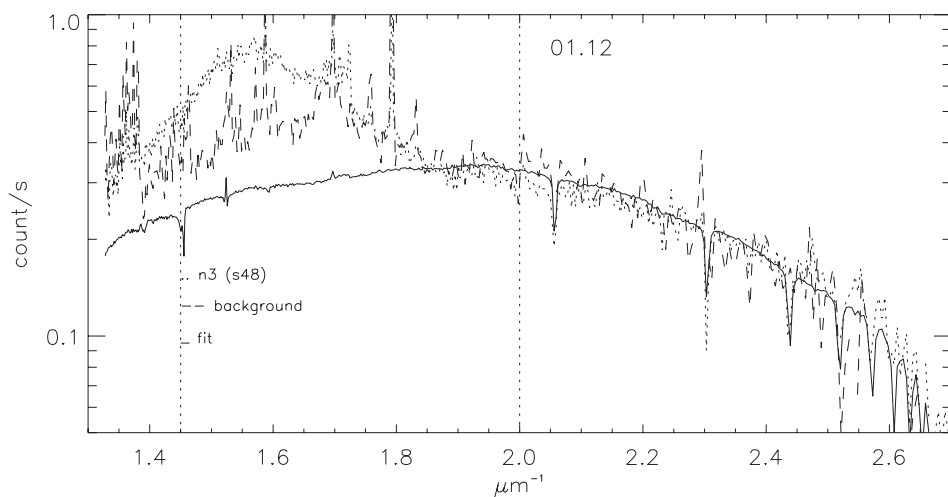


Figure 11. Position 3 main spectrum (s48, dots), the background close to the nebula in the same observation (times 5.5, solid line), and the fit, proportional to the spectrum of HD 44179 (first observation) and to $1/\lambda$: the blue part of the nebular spectra $n3$, $n4$, $n5$, is light from HD 44179 scattered in the atmosphere.

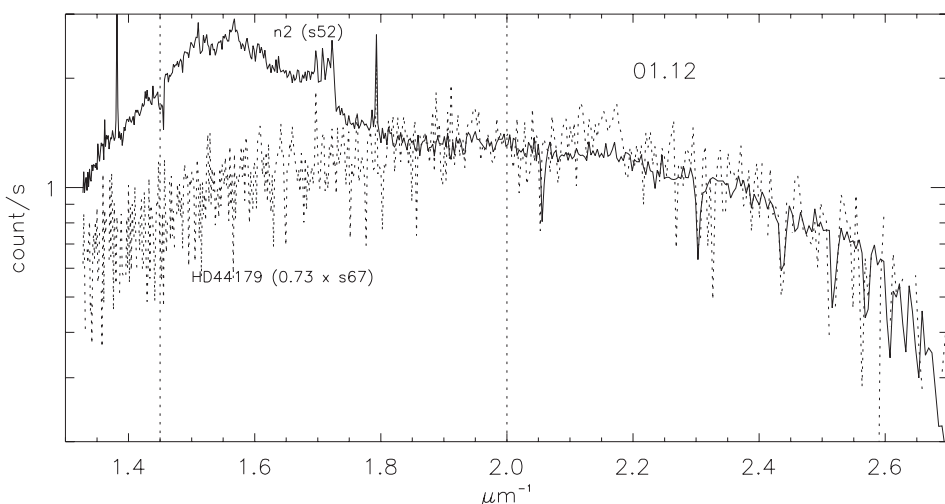


Figure 12. The main spectrum, s52, from $n2$ 2-D array, is observed at about the same distance from HD 44179 as s67 and s68, from the first observation of HD 44179, are. The blue part of the main spectrum s52 at position $n2$ is in-between the blue parts of s67 and s68. As shown by the plot, it is perfectly reproduced by one of these spectra with an appropriate rescaling (here s67 of HD 44179's first observation times 0.73).

distance from HD 44179 as s52 (of $n2$) is. The blue parts of the other spectra of $n2$ 2-D array are either like that of s52 (for pixel numbers close to 52), or, like the blue part found previously at positions $n3$, $n4$, $n5$.

It will be concluded that for all spectra in the nebula, only the red part of the spectra can safely be attributed to light from the nebula. In the blue, the spectra are dominated

either by direct, refracted, light from HD 44179, or by light from HD 44179 forward scattered by atmospheric aerosols.

One can wonder why $s52$ of $n2$ is three times larger than $s48$ of $n5$ (Fig. 10), while position $n5$ is closer to HD 44179 than $n2$ is. The reason might be the increase of airmass between the two observations (Table 1), which diminishes the amount of direct light from HD 44179 and increases the scattered light in the $n5$ observation.

4. February 2002 observations

There are three February 2002 observations, one of HD 44179, one $5.4''$ north of HD 44179 (position 1), and one of HD 44113.

The two star observations are plotted in Fig. 13. The decrease of the signal, when moving away from the main pixel ($s35$ and $s36$ respectively for HD 44179 and HD 44113, solid line on the plots), is similar in both observations, although it is faster for HD 44179 (while HD 44179 and HD 44113 have comparable magnitudes): it occurs within 1–2 pixels ($\sim 4''$) from $s35$ for HD 44179, within 4 pixels ($\sim 8''$) for HD 44113 (as for the December 2001 observation of HD 44179). This may be related (Zagury 2005) to the exposure time, which, for HD 44179, has been reduced from 15 seconds in December 2001 to 5 seconds in February 2002. The exposure time for HD 44113 is 15 seconds, as it was for the December 2001 observations of HD 44179.

The spectrum of the nebula clearly appears in the red, superposed on the spectrum of HD 44179, for spectra $s33$ and $s32$, $s37$ and $s38$. The nebular contribution to the spectra is far more significant in this observation than it was for the December 2001 ones, which, *a priori*, can be due either to a slightly different position of the slit axis ($3''$ wide) in declination, or, more likely, to a difference in atmospheric extinction between the two nights (which implies a lower extinction for February 2002 observations and could be related either to the lower airmass (Table 1), or, to the smaller exposure time of these observations). The nebular spectrum is present in the red only and does not seem to perturbate the spectrum of HD 44179 in the blue.

The nebular 2-D array was presented, along with the one of December 2001 observation at the same declination, in Fig. 5. Two main spectra ($s32$ and $s33$) merge out from the 2-D array (rather than one in December 2001 observations), which is to be attributed to the binning by four procedures used in the data reduction of February observations. The shape of $s33$ (a fit is proposed in section 6.), and the decrease of the spectra on each side of the central pixels, indicate, as for $n1$ December 2001 observation, that the spectra are dominated in the blue by direct light from HD 44179: the nebular spectra are superposed on HD 44179's spectra, and restricted to the red. The background reaches the night sky spectrum at the edges of the slit.

5. March 2003 observations

5.1 HD 44179

The two March 2003 observations of HD 44179 are similar, and were discussed in Zagury (2005) (Figs. 1 and 2). As for the February 2002 observation, direct light from HD 44179 is spread over $\sim 8''$, i.e., $\sim 4''$ (3–4 pixels) on each side of the central pixel. The noise level is ~ 1 count/s, higher than in December 2001, which is due to the smaller exposure time.

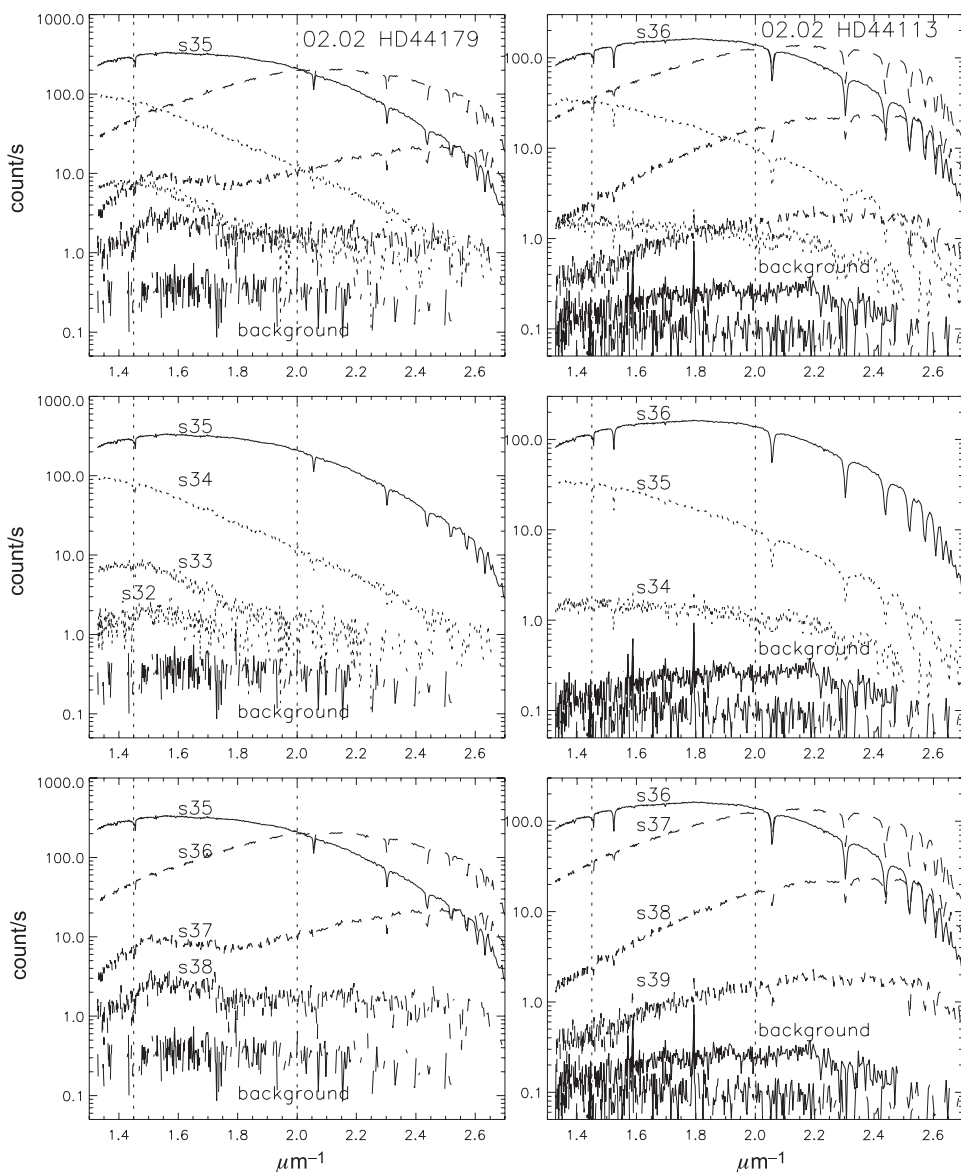


Figure 13. February 2002 observations of HD 44179 (left column) and HD 44113 (right column). The spectrum of the nebula appears in the red, above the spectrum of HD 44179, for spectra *s37*, *s38*, and *s33*. Two backgrounds are plotted for HD 44113, to illustrate its decrease with distance from the star.

5.2 Nebula

The March 2003 observations of the Red Rectangle nebula are classified in Fig. 14, from left to right and top to bottom, by decreasing importance of the signal. The main features are the same as for the December 2001 observations.

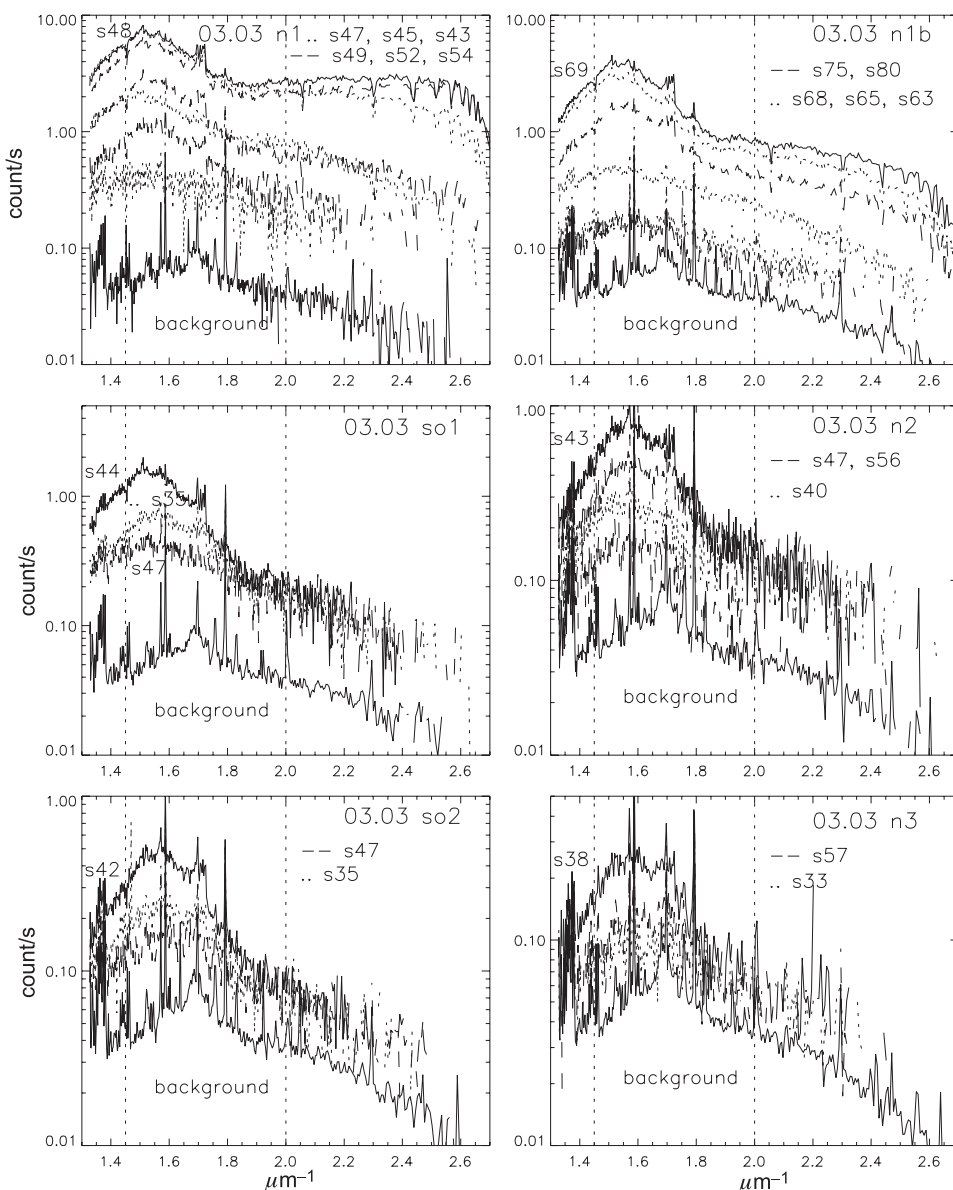


Figure 14. March 2003 observations of the Red Rectangle nebula. Plots are ordered from left to right and top to bottom by decreasing importance of the signal (i.e., by increasing distance from HD 44179). For the sake of clarity, few spectra only, for each observation, are represented.

The backgrounds away from the nebula are identical (Fig. 15, left), in shape and intensity. The match with Massey & Foltz (2000) night sky spectrum is good except in the blue. The excess found in the blue reproduces the background of HD 44179 observations; it is proportional to the product of HD 44179 1-D spectrum (first observation) by $1/\lambda$, and is interpreted as remaining light from HD 44179 scattered in the atmosphere (Fig. 15, right).

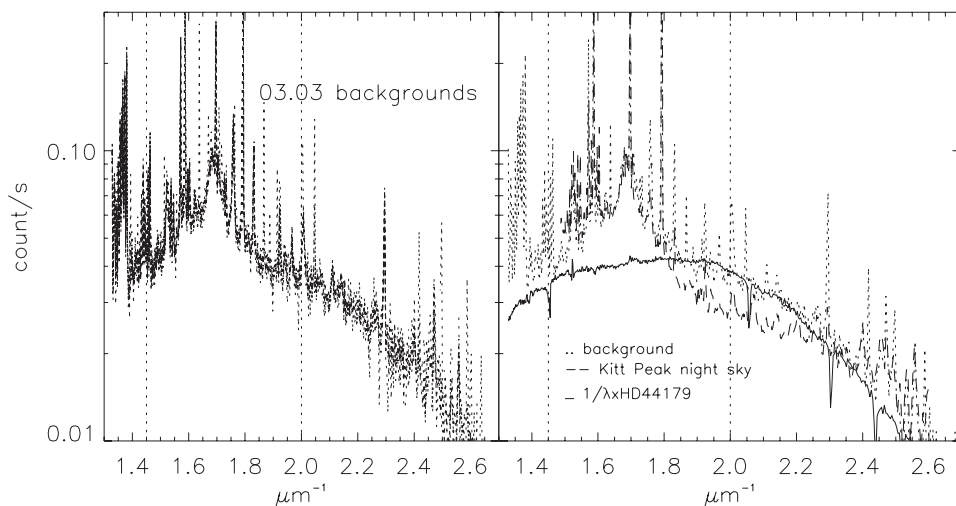


Figure 15. Left: The backgrounds from Fig. 14 (March 2003 observations of the nebula) superimpose well, as for December 2001. Right: The night sky spectrum of Massey & Foltz (2000) (dashes) is compared to the average of the background spectra (dots). In the blue the average spectrum presents an excess, due to light from HD 44179 scattered in the atmosphere, fitted (solid line) by the spectrum of HD 44179 multiplied by $1/\lambda$ (and a constant factor).

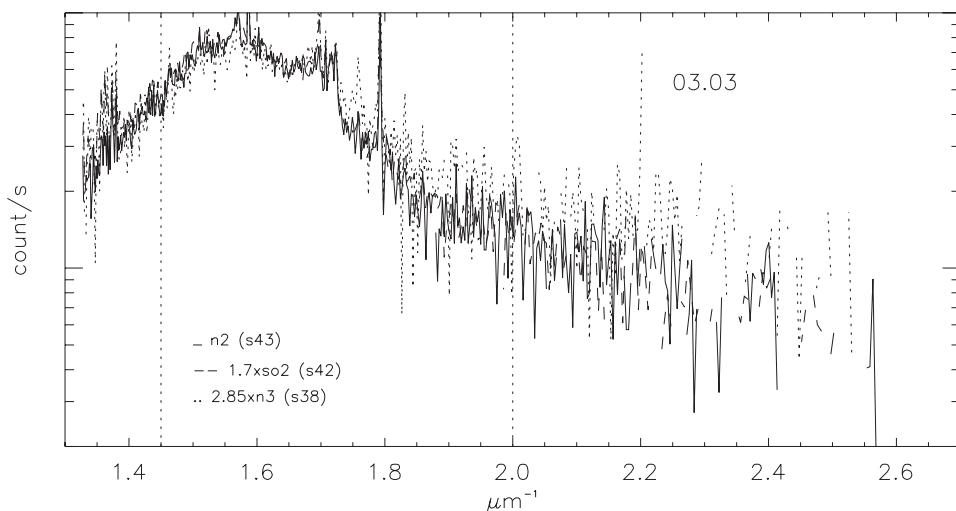


Figure 16. The spectra at positions $n2$, $n3$, $so2$ of March 2003 observations are identical, up to a constant factor ($n2 = 2.85$, $n3 = 1.7 so2$).

The main spectra at positions $n2$, $n3$, and $so2$ are identical, up to a constant factor (Fig. 16).

In the red (Fig. 14), the red bump of the Red Rectangle is present, within the nebula, in all observations. The red rise keeps a constant slope around the main pixel of an observation, and will diminish toward the edges of the nebula; the bump disappears when the pixel is out of the nebula.

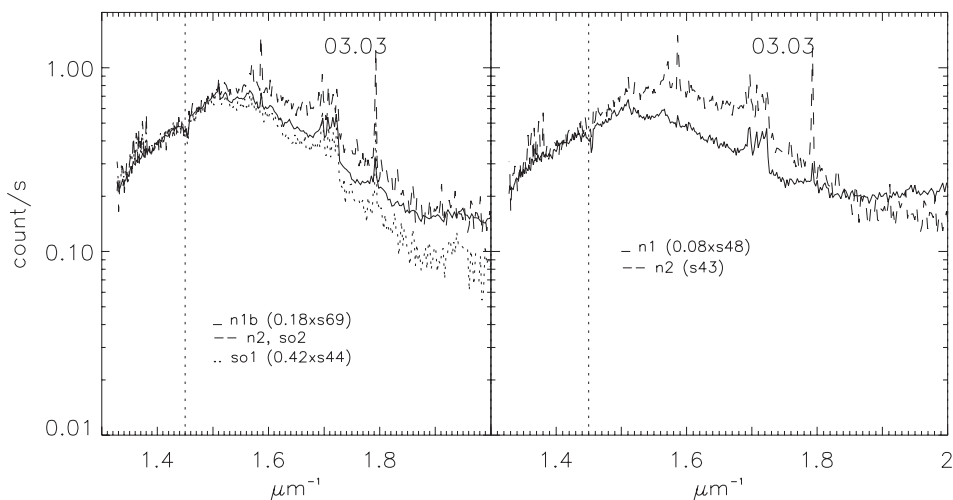


Figure 17. Left: The red slopes of all nebular observations (except *n1*) are identical. Here, the spectrum of the nebula averaged over positions *n2* and *so2*, dashes, is compared, in the red, to the main spectra at positions *n1b* (solid line) and *so1* (dots). Right: The red rise of the main spectrum *s48* at *n1* is identical to the precedings in its reddest part only. From the ozone absorption region onwards it is less steep.

The red rises of the main spectra of all observations (except *n1*) are proportional (Fig. 17, left). *n1* main spectrum shares the same slope in the reddest part of the red rise, but is less steep from the ozone absorption region onwards (Fig. 17, right).

The shape of the blue spectra, for a given 2-D array of one of the four observations with lower signal (*n2*, *n3*, *so1*, *so2*), is the same in and out of the nebula, down to the background (Fig. 14). All four observations blue spectra are proportional (Figs. 16 and 18). One will deduce that, as for December 2001 observations, the blue part of these spectra is mainly light from HD 44179 scattered in the atmosphere (Figs. 16 and 18); and that the difference between the spectra at *so1* and at positions *n2*, *n3*, *so2*, is due to the relative proportion in the spectra of light from HD 44179 scattered in the atmosphere.

The blue decrease of the main spectrum (*s69*) at position *n1b* perfectly reproduces the blue part of spectrum *s54* (scaled by a factor of 0.1) of the first observation of HD 44179 (Fig. 19).

n1 (Fig. 14) is the nebular observation with the highest signal and the closest to HD 44179. The nebula bump is perceived from pixel 43 to pixel 56. The main spectrum is *s48*, which is flat in the blue. Such a flat blue slope is observed in the February 2002 observation of HD 44179 (*s38*, Fig. 13, for instance), and can be attributed to direct light from HD 44179. Spectra close to *s48* differ by a gas extinction factor only ($\propto e^{-g/\lambda^4}$, Fig. 20), as it was shown to be the case with the spectra from HD 44179's 2-D array (Zagury 2005). Away from pixel 48, the spectra match that of other observations in the nebula, observed at comparable distances from HD 44179: spectrum *s52* of *n1* for instance is identical (with a scaling factor of 1.3) to the main spectrum, *s69*, of *n1b*, the blue part of which was shown to be proportional to *s54* from the first observation of HD 44179 (Fig. 19).

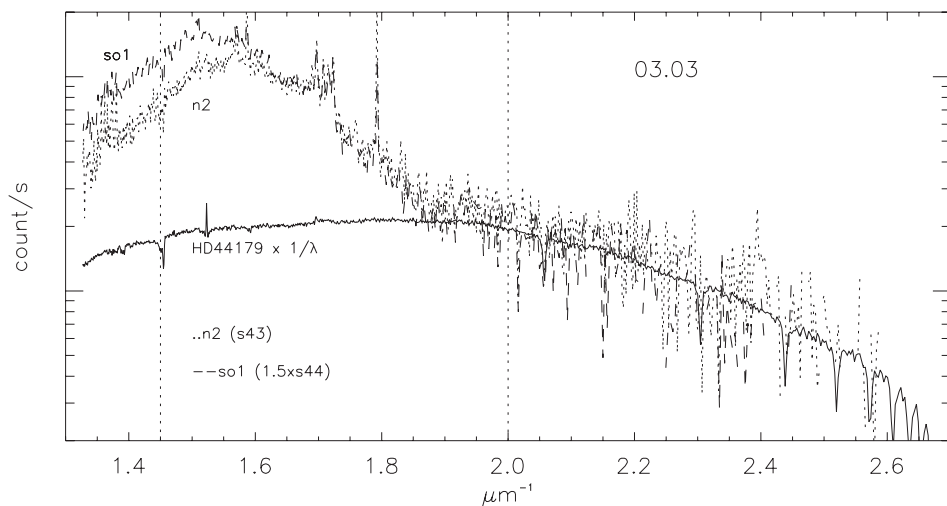


Figure 18. Main spectra at positions *n2* (*s43*) and *so1* (*s44*) are proportional in the red (Fig. 17), but also in the blue. Their blue part reproduces the background of HD 44179 observations; it is here fitted by the spectrum of light from HD 44179 scattered in the atmosphere (proportional to $1/\lambda$ and to the spectrum of HD 44179).

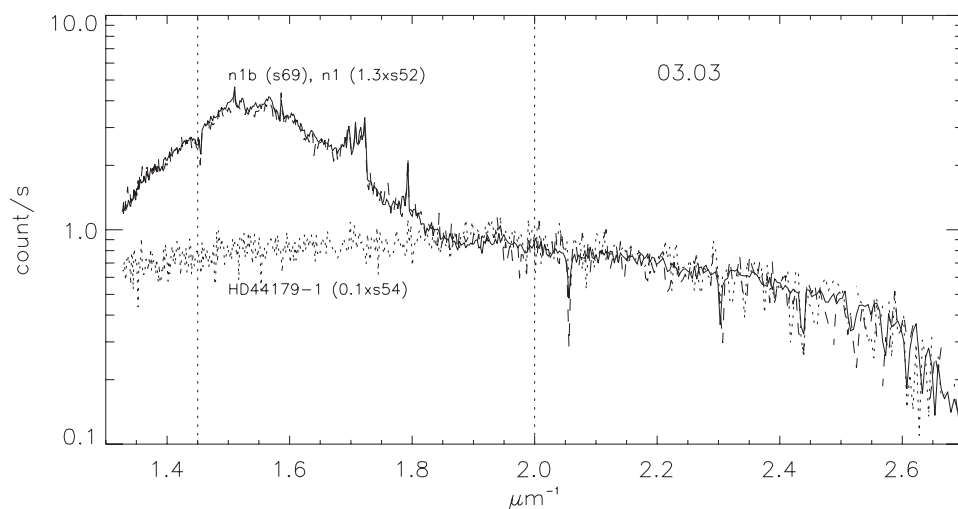


Figure 19. The blue part of March 2003 observation *n1b* (*s69*) is reproduced by spectrum *s54* of the observation of HD 44179 (with a decrease of 10 in intensity): close to the star, the spectrum of the nebula appears over the spectrum of refracted-scattered (in the atmosphere) light from HD 44179. The same conclusion was reached from the February observations of HD 44179 (see Fig. 13). Spectrum *s52* (dashes, scaled by a factor 1.3), from *n1* observation, superimposes well on *n1b*'s *s69*.

As for December 2001 or February 2002, the spectra of the nebula observed in March 2003 reveal the light from the nebula in their reddest part (where atmospheric extinction is minimized, before ozone absorption), but are, in the blue (after the ozone

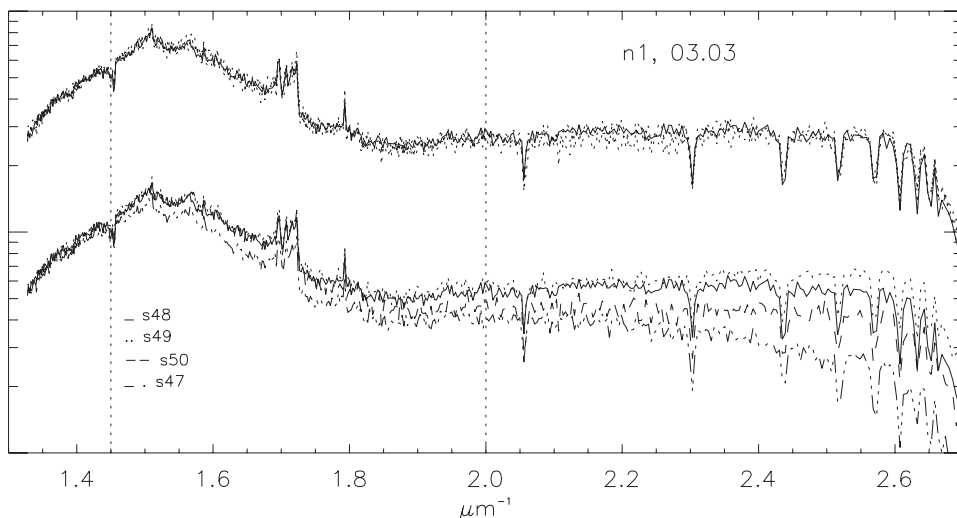


Figure 20. Bottom spectra: s47, s48, s49, s50 from n1 March 2003 observation, scaled to the same value in the red. Top spectra: The blue part of the spectra differ by their blue slope only, which is determined by the Rayleigh extinction in the atmosphere. s48 is well reproduced by $0.9e^{0.018/\lambda^4}$ s47, $1.3e^{-0.004/\lambda^4}$ s49, $1.6e^{0.006/\lambda^4}$ s50.

absorption region), dominated by light from HD 44179 – direct light or light scattered in the earth atmosphere.

6. Comparison of the observations at different dates

Figure 21 compares the backgrounds, observed away from HD 44179, of the three observing runs. In the red, a scaling factor close to 1 will give a good superimposition of the spectra. In the blue, the three backgrounds differ by the amount of light from HD 44179 scattered in the atmosphere. The difference between December 2001 and March 2003 backgrounds is small. It becomes significant with the February 2002 background, which is understandable since only the closest position to HD 44179 was observed in February 2002. We conclude that the night sky was nearly the same for the three runs. It implies that the conditions of observation, the sensitivity of the detector and the atmospheric extinction are comparable.

The 1-D spectra of HD 44179 observed at different dates superimpose well after correction for a difference of Rayleigh extinction (Fig. 22) between the observations.

Main spectra at positions n2 and n3 are proportional in March 2003 (section 5.), while they have different shapes in December 2001 observations (section 3.). The pointing of the telescope could be responsible for part of this discrepancy, but the most reasonable explanation is that the difference of optical depths (airmass) between the two dates changes the proportion of refracted and scattered light, from HD 44179, in the spectra. It means that the standard way of reducing data, subtraction of a background and normalization by the spectrum of a reference star, cannot be applied to these observations.

A precise position-to-position comparison of December 2001 and March 2003 observations (Fig. 23), for positions n2 and n3, shows that the red rise of the nebular

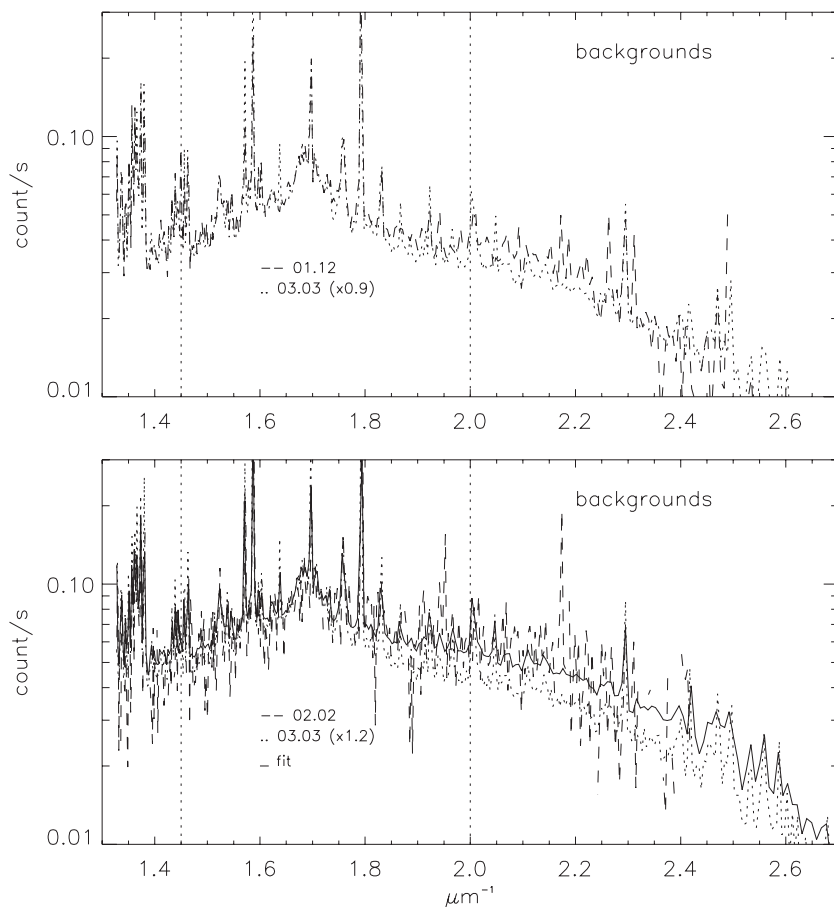


Figure 21. Comparison of the backgrounds. Upper plot: December 2001 and March 2003 backgrounds are identical, except for a small difference in the blue. Lower plot: February 2002 and March 2003 backgrounds differ in the blue, because of the importance of the scattered light (from HD 44179) in the atmosphere in the February background. February 2002 background is well fitted by the sum of March 2003 background, and a component of light from HD 44179 scattered by the atmosphere (proportional to the February 1-D spectrum of HD 44179 and to $1/\lambda$).

spectra is identical from one date to the other, while the spectra differ in the blue from the ozone absorption region onwards.

$n1$ is the nearest position to HD 44179 observed in the nebula, and the one for which direct light from HD 44179, refracted in the atmosphere, enhances the level of the spectrum far above that at the other positions. It is also the position for which differences between the different dates are the most striking (Fig. 24). The exact position of the slit axis with respect to HD 44179 can influence the level of the spectrum. The high airmass ($A.M = 1.51$, against 1.41 for December 2001 and 1.38 for February 2002) probably contributes to the low level of the March 2003 spectrum, and its steep rise in the red, which is close to what is observed at larger distances from HD 44179: an increase of extinction diminishes the direct light from HD 44179, which is the main

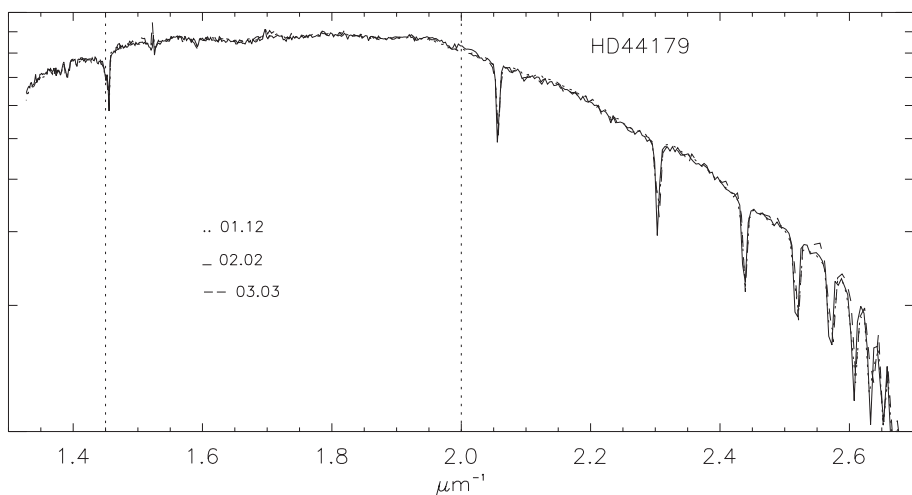


Figure 22. The 1-D spectra of HD 44179 for the three runs differ by an atmospheric transformation: 1-D spectra of the first observations of December 2001 and March 2003 superimpose on the February 2002 spectrum after multiplication by $1.24e^{0.008/\lambda^4}$ (for December 01), and $1.5e^{0.02/\lambda^4}$ (for March 03).

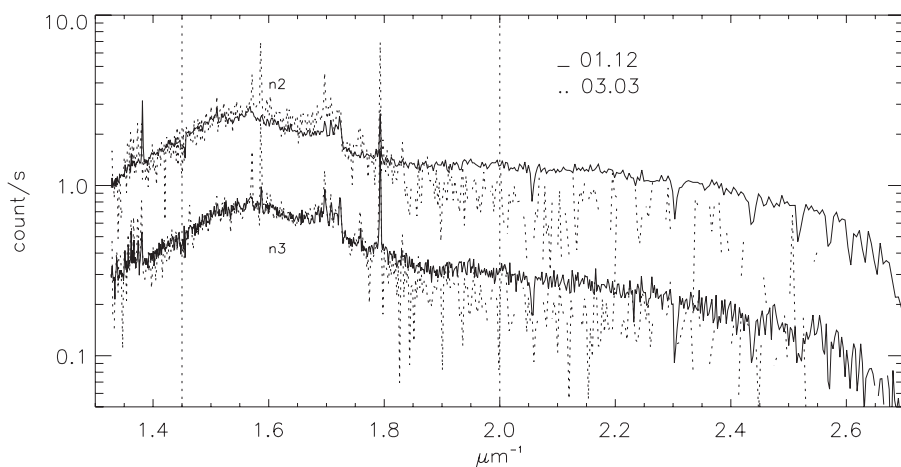


Figure 23. Spectra observed at $n2$ and $n3$ on different dates are proportional in the red but differ by their blue continuum. $n2$ and $n3$ spectra of March 2003 have been multiplied by 7 and 3 to match December 2001 observations.

component of the $n1$ spectra in the blue, and the reason for its high level; the Red Rectangle nebular bump is then highlighted.

This further raises the question of the influence of the blue continuum on the slope of the red rise at $n1$, for December 2001 and February 2002 observations. Figure 24 (right plot), for instance, shows that it is possible to reproduce February 2002 $n1$ spectrum by the sum of spectra extracted from the 2-D arrays of December 2001 observations of HD 44179 and of position $n2$. We also note, still concerning the 2-D array of February 2002 observation, that when the blue continuum diminishes, the slope of the red rise

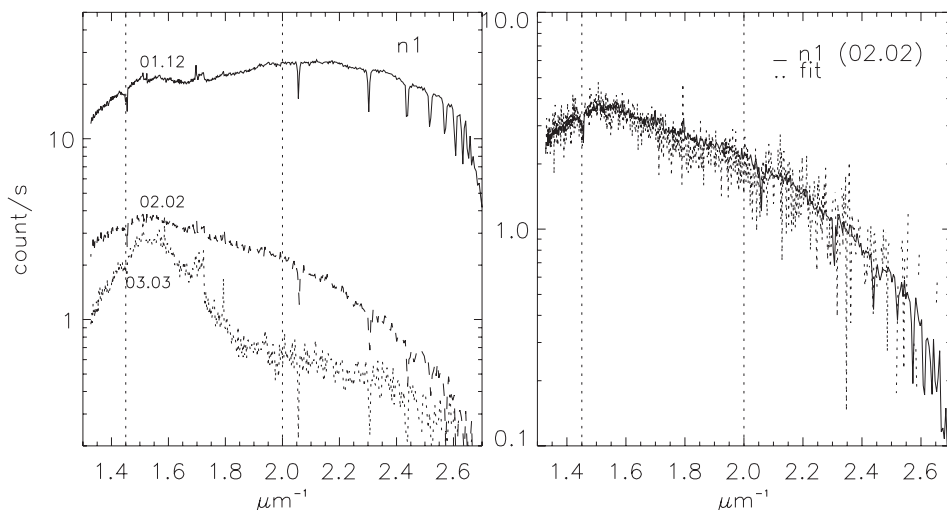


Figure 24. Main spectra at $n1$ are extremely different from one date to the other (left). The difference is explained by the position of the slit and the airmass, which determines the proportion of refracted and scattered light (in the atmosphere) from HD 44179 in the spectrum. The right plot shows that the main spectrum of February 2001 $n1$ observation can be reproduced by a combination of spectra from HD 44179, and from the nebula at position $n2$, from December 2002 observation.

increases to its value at positions farther away from HD 44179 ($s28$ of $n1$, February 2002 observation, for instance, is identical to the main spectrum $s49$ of $n2$, December 2001 run).

7. Discussion

The previous analysis leads to two conclusions. The first is that the only part of the spectra which can be attributed to light from the nebula is the red bump, the spectrum at smaller wavelengths being light from HD 44179 refracted or scattered in the atmosphere. The second conclusion is that only the red rise of the bump is not modified from one observation to another.

What is the meaning of the bump, and where is the light from the central star scattered by the nebula? Either the light from HD 44179 scattered by the nebula really is reduced to a small offset, in the blue and maybe in the red; or, it is seen in the red only because it is strongly extinguished from the ozone absorption region onwards (and towards the UV).

In the first case, the red bump could be, as it has been admitted so far, an emission process due to some kind of particle which would absorb light in the UV and fluoresce in the red. Light from HD 44179 scattered by the nebula is then negligible compared to the fluorescence process. However, no such particles have ever been synthesized on earth with the required spectrum, and their nature is today, after some thirty years of research, still a speculative subject (Li & Draine 2002; Van Winckel *et al.* 2002).

Also in this case, the red bump could result from extinction and scattering by large grains or by the gas (such an explanation was suggested by Cohen *et al.* (1975), see

section A (Appendix)). Forward scattering by large grains would give a spectrum $\propto e^{-a/\lambda}/\lambda$ (a a constant), with a much too large bump to fit the data (all spectra of the Red Rectangle nebula found in the literature give a ratio of FWHM to central wavelength in the range 0.18–0.30, while scattering by large grains would imply a ratio close to 4). It is also not compatible with the large angles of scattering involved in the present case. Scattering by gas will have a spectrum $\propto e^{-g/\lambda^4}/\lambda^4$ (Rayleigh 1871), which implies a constant ratio of FWHM to central wavelength of 0.66, also too large to match the observations.

In addition to the absence of scattered light from the nebula in the blue, if the red bump is a true nebular feature (an emission process or due to scattering), it is still necessary to understand the reasons for the variability of the blue decrease of the bump (Figs. 10, 17, 20, 23), while the red rise keeps, in general, a constant slope (modifications of the red slope are observed close to HD 44179 only, and may be, as pointed out in section 6., due to the underlying continuum of HD 44179 light, refracted or scattered in the atmosphere). The correlation between the modifications of the blue side of the bump, and the ozone absorption region, would have to be pure coincidence, as would have to be the transformations used to superimpose the main spectra of *n*1 March 2003 observation (Fig. 20), which attest the role of atmospheric extinction on the decrease.

On the other hand, atmospheric extinction would well explain the particularities of the Red Rectangle nebula spectra. The red rise of the bump is constant, within the nebula, in all observations. It corresponds to a wavelength region with no ozone absorption and where Rayleigh extinction is low. The progressive interruption of the red rise, and the blue decrease, both coincide with the wavelength region where ozone absorption operates. At larger wavelengths, Rayleigh extinction by nitrogen ($\propto e^{-g/\lambda^4}$) will relay ozone absorption.

If atmospheric extinction is the explanation of the red bump, the true meaning would be that atmospheric extinction does not act in the same way for the nebular light as it does for point sources' light, for HD 44179 or HD 44113 for instance. More generally, it implies that extended objects are not extinguished by the atmosphere in the same way as point sources are, atmospheric extinction being far more efficient on the former. This can only be related to turbulence, and to the structure of the atmosphere, already responsible for such phenomena as scintillation, which affects starlight and not the light from planets.

This hypothesis is further supported by the observations presented in Zagury (2005). Faint and extended galaxies observed with FAST (see the FAST database at: <http://tdc-www.harvard.edu/archive/>), as UGC11917, also show a strong decrease from the ozone absorption region onwards, and their blue shape is close to that of the night sky. Except for the three nebulae used in Zagury (2005), I didn't have access to the 2-D spectra of the Kwitter–Henry database, but anyone who considers the continua of the Kwitter–Henry spectra will observe that they can be classified, according to their shape, into one of very few categories (I found 4 well-defined categories, and this classification can probably be simplified). Many of these spectra have the same marked feature in the red as the Red Rectangle has, while others have a star-like spectrum. One can surmise that this classification results from the relative proportion, in the spectra, of light from the central star refracted by the atmosphere, scattered light from the central star in the atmosphere, night sky, and nebular light in the red.

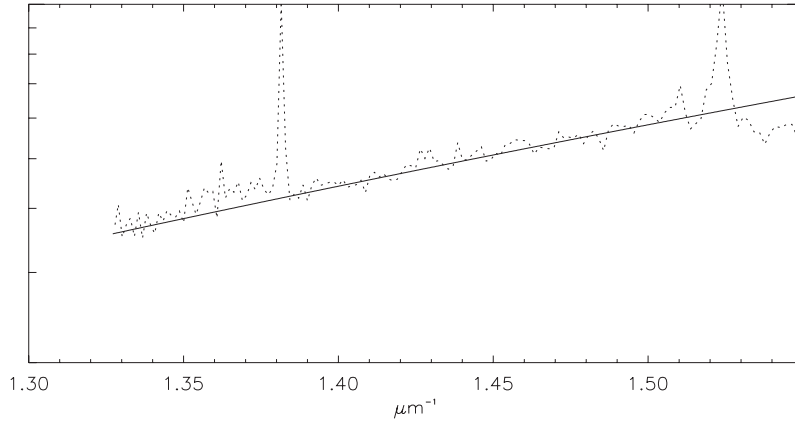


Figure 25. In dots, the red slope of the Red Rectangle nebula normalized by the spectrum of the A0 star HD 44113 (the y-axis is logarithmic, the scaling is arbitrary). It behaves as $1/\lambda^4$ (solid line), indicating scattering by the gas.

The only part of the Red Rectangle nebula spectrum which is reliable, because of the small effect of atmospheric extinction in this wavelength region, is the red part, under $\sim 1.5 \mu\text{m}^{-1}$ (above $\sim 6600 \text{ \AA}$), before ozone absorption modifies the spectrum. To investigate the nature of this part of the spectrum, we need to study its power law dependence when divided by the spectrum of an A0 star (same spectral type as HD 44179, $T \sim 7500^\circ\text{K}$), as HD 44113. Figure 25 shows that the red rise of the nebula spectrum has a $1/\lambda^4$ dependence, which indicates Rayleigh scattering by the gas. This would be in conformity with the conditions which prevail for these observations, i.e., a scattering medium close to the source of light, and large angles of scattering.

The conclusions finally adopted here (that atmospheric extinction is more effective on extended objects than it is on point sources, and that the true spectrum of the Red Rectangle nebula reflects scattering by the gas (hydrogen)) can easily be checked by observing the Red Rectangle nebula, and a non-reddened star of A0 type, as HD 44113, from space, on a large wavelength range. In the absence of atmospheric extinction, we expect the ratio of the spectra to vary as $1/\lambda^4 e^{-g/\lambda^4}$ over all the visible and UV wavelength range. It would also be desirable to observe different positions in the nebula, close to, and far from, HD 44179, to determine the variations of the intensity of the scattering, and to see if, close to the star, scattering by large grains, with a $1/\lambda$ dependence, is detected.

8. Conclusion

The observations of the Red Rectangle nebula presented in this paper outline the importance of atmospheric extinction on the re-distribution of starlight in the vicinity ($\sim 1'$) of a star. The spectra of the nebula all appeared to be the superimposition of two components, one in the red, clearly associated with the nebula and one in the blue, which is mainly either direct light from HD 44179 refracted by the atmosphere, or light from HD 44179 forward scattered by atmospheric aerosols (the proportion of direct and scattered lights depends on the angular distance from HD 44179, and on the degree of atmospheric extinction).

Within the precision of these observations, in none of the spectra, observed at different places of the nebula, and at different dates, was it possible to detect the spectrum of the nebula in the blue, which means that either there is no (or a negligible) nebular continuum in the blue, or, that atmospheric extinction strongly extinguishes the nebular light from the ozone absorption region on, towards the UV.

The correlation between the decrease (the blue side of the red bump) of the light from the nebula and the ozone absorption region, and the variations of the spectra at one position from one date to another, naturally favours the second hypothesis. It will then be inferred that faint extended objects are more sensitive to atmospheric extinction than point sources are. Hints to the same conclusion can be found in the spectra of the Kwitter–Henry database, and of faint extended galaxies observed with FAST.

This is the first and main conclusion of this paper. It implies that it is not possible to obtain a reliable spectrum of faint extended objects, covering the whole visible spectrum and in ground-based observations, by the usual way of normalizing the spectrum by the spectrum of a reference star.

These considerations give an alternative solution to the Red Rectangle problem, and more generally to the red bump found in the spectrum of some nebulae. To explain the bump, the last thirty years of research have focused on the existence of particles which would absorb UV light and fluoresce in the red (section A). Such particles, which should be abundant enough to account for some 30% of the scattered red light in the galaxy (Szomoru & Guhathakurta 1998), with an intrinsic photon conversion efficiency of the photoluminescence near 100% (Zubko *et al.* 1999; Van Winckel *et al.* 2002), would indeed represent a remarkable discovery, and it is a puzzling question to understand how particles we are not able to synthesize on Earth, can be so abundant in the cold and nearly empty interstellar medium. However, none of the attempts to identify these particles has given a satisfactory result yet, their existence being as speculative today as it was thirty years ago (Van Winckel *et al.* 2002).

That atmospheric extinction be responsible for the red bump would certainly lead to a review of the problem on a more simple and pragmatic basis. It will also explain the similarity outlined by Zagury & Fujii (2003) between the spectrum of a red horizon at sunrise and the spectra of red nebulae. Both spectra result from a combination of Rayleigh scattering by a gas (nitrogen for the red horizon, hydrogen for the nebula), and of atmospheric extinction (Rayleigh extinction and ozone absorption). This is the second conclusion of the paper.

One way to verify these two conclusions would be to obtain broad-band spectra of extended objects from space, with the Hubble Space Telescope for instance. Concerning the Red Rectangle, I expect the ratio of spectra of the nebula and of a non-reddened A0 star, as HD 44113, to give a $1/\lambda^4$ dependence over the complete visible spectrum. To observe Rayleigh scattering by the gas, observed positions in the nebula should be taken far enough from HD 44179. Close to HD 44179 it is possible that forward scattering by large grains will be more efficient than scattering by the gas, and that a $1/\lambda$ dependence be found. These observations will evidence, for the first time, Rayleigh scattering by the gas in the interstellar medium.

A refined study of the spectra from the Kwitter–Henry database, at the early stage of data reduction where atmospheric extinction is not yet corrected, could also provide valuable information on the relationship between the conditions of observation and the shape of the spectra. As already mentioned in Zagury (2005), there is little doubt

that the proximity of the central star, the balance between the proportions of light from the central star refracted or scattered in the atmosphere, and from the night sky, have a determining influence on the final shape of the spectra in the blue. For nebulae with a reliable spectrum in the red (where atmospheric extinction is low), the ratio of scattered light from the nebula to the spectrum of an unreddened star of the same spectral type as the central star, should also provide informations on the type of particles which scatter starlight (large grains with a $1/\lambda$ dependence, hydrogen for Rayleigh scattering by the gas), and thus on the geometrical relationship (proximity, angles of scattering) between the nebula and the star.

A difference in the effect of atmospheric extinction on point sources and on extended objects could be due to the time-dependent structure turbulence introduces in the atmosphere. The consequences for our understanding of the atmosphere, like the relations which exist between the spectra at very short angular distances from a star, would merit further attention and investigation.

Acknowledgements

This work was made possible thanks to the courtesy of Roc Cutri and John Huchra. Most of all, I am indebted to Lucas Macri, for his unfailing kindness in repeating the observations of the Red Rectangle.

Appendix

The Red Rectangle

The large interest (Simbad quotes some 320 references to the Red Rectangle) raised in the thirty past years by the Red Rectangle takes its origin in the Cohen *et al.* (1975) article, ‘The peculiar object HD 44179 (“The Red Rectangle”)’. In conclusion of their paper, the authors address several problems which concern the star system at the center of the nebula, the nature of the bi-conical nebula, and the reason(s) for its important brightness in the red.

Large advances in the comprehension of the inner part of the nebula have been made in the 1990s, from high resolution imaging, with Roddier *et al.* (1995) who showed that we do not see direct starlight from HD 44179, only scattered light escaping from the poles of an optically thick disc, seen edge-on. The disc, more likely a torus (Tuthill *et al.* 2002), surrounds a spectroscopic binary with an orbital period of 300 days (Van Winckel, Waelkens & Waters 1995). The optical light escaping the torus is that of an A0 star, with effective temperature $\sim 7500^\circ\text{K}$ (Waelkens *et al.* 1996). A schematical representation of this inner system (~ 100 mas around HD 44179) can be seen in Fig. 3 of Waelkens *et al.* (1996) or Fig. 8 of Men’shchikov *et al.* (1998). Recent high resolution images (Tuthill *et al.* 2002; Cohen *et al.* 2004) of this inner part of the nebula seem to confirm this representation.

On a larger scale the nebula is biconical, and occupies $\sim 40'' \times 40''$ on the sky. The question of its red colour, addressed by Cohen *et al.* (1975), is still a subject of debate. The quite logical proposal by Cohen *et al.* (1975), that it could be light from the central A0 star extinguished (by small and large grains) and scattered by the gas (because of the large angles of scattering), did not attract much attention (except for two papers, Greenstein & Oke 1977; Perkins *et al.* 1981), on-going research being

oriented towards the search of particles which would absorb in the UV and luminesce in the red (Duley 1985; d'Hendecourt *et al.* 1986; Seahra & Duley 1999). However, the existence of these particles is still in the speculative domain (Li & Draine 2002; Van Winckel *et al.* 2002).

References

- Cohen, M. *et al.* 1975, *ApJ*, **196**, 179.
Cohen, M., Van Winckel, H., Bond, H. E., Gull, T. R. 2004, *Astron. J.*, **127**, 2362.
Duley, W. W. 1985, *MNRAS*, **215**, 259.
Fabricant, D. *et al.* 1998, *PASP*, **110**, 79.
Greenstein, J. L., Oke, J. B. 1977, *PASP*, **89**, 131.
d'Hendecourt, L. B., Léger, A., Olofsson, G., Schmidt, W. 1986, *A&A*, **170**, 91.
Li, A., Draine, B. T. 2002, *ApJ*, **564**, 803.
Massey, P., Foltz, C. B. 2000, *PASP*, **112**, 566.
Men'shchikov, A. B., Balega, Y. Y., Osterbart, R., Weigelt, G. 1998, *NewA*, **3**, 601.
Perkins, H. G., Scarrott, S. M., Murdin, P., Bingham, R. G. 1981, *MNRAS*, **196**, 635.
Lord Rayleigh 1871, *Phil. Mag. XLI*, **274**, 107.
Roddier, F., Roddier, C., Graves, J. E., Northcott, M. J. 1995, *ApJ*, **443**, 249.
Seahra, S. S., Duley, W. W. 1999, *ApJ*, **520**, 719.
Szomoru, A., Guhathakurta, P. 1998, *ApJ*, **494**, L93.
Tokarz, S. P., Roll, J. 1997, *A.S.P. Conference Series*, Gareth Hunt and H. E. Payne (eds.) **125**, 140.
Tuthill, P. *et al.* 2002, *A&A*, **389**, 889.
Van Winckel, H., Waelkens, C., Waters, L. B. F. M. 1995, *A&A*, **293**, L25.
Van Winckel, H., Cohen, M., Gull, T. R. 2002, *A&A*, **390**, L47.
Waelkens, C., Van Winckel, H., Waters, L. B. F. M., Bakker, E. J. 1996, *A&A*, **314**, L17.
Zagury, F. 2005, 'Spectral analysis of the background in ground-based, long-slit spectroscopy', this issue.
Zagury, F., Fuji, M. 2003, *NewA*, **8**, 549.
Zubko, V. G., Smith, T. L., Witt, A. N. 1999, *ApJ*, **511**, L57.

Frozen Hydrocarbon Particles of Cometary Halos as Carriers of Unidentified Emissions

Irakli Simonia

Georgian National Astrophysical Observatory, Kazbegi Av., 2a, Tbilisi 0160, Georgia.

e-mail: irsim@geo.net.ge

Received 2004 June 9; accepted 2005 July 13

Abstract. The possible nature of unidentified cometary emissions is under discussion. We propose a new model of the ice particles in cometary halos as a mixture of frozen polycyclic aromatic hydrocarbons and acyclic hydrocarbons. We describe principal properties of frozen hydrocarbon particles (FHPs) and suggest interpreting some of the unidentified cometary emission lines as the photoluminescence of FHPs. The results of comparative analysis are present.

Key words. Comet—ice—luminescence—frozen hydrocarbon particles—unidentified lines.

1. Introduction

The problem of unidentified cometary emissions in optical part of a cometary spectrum is known to remain of great interest. Such emissions are unidentifiable both numerically and comparatively, and are usually presented in the form of tables or data base (Brown *et al.* 1996; Cochran & Cochran 2002). Their identification depends on the methods used to obtain and process the spectral data, the amount of accumulated comparative laboratory results, knowledge of other excitation mechanisms for the corresponding emission lines. Although unidentified emissions may emerge within an ordinary resonance–fluorescence mechanism due to transition of corresponding atoms and molecules from higher levels, there still remains a possibility of photoluminescence excited in a solid cometary substance exposed to solar UV radiation (Simonia 1999). This process appears to be highly probable, particularly, if we take into account considerable amount of organic components in the cometary ice. Polycyclic Aromatic Hydrocarbons (PAHs) belong to a most important class of chemical compounds found in cosmic bodies. There is no doubt that they are contained in the cometary substance (Ehrenfreund *et al.* 2000). We suggest that ice of a cometary nucleus may contain mixtures of PAHs and acyclic hydrocarbons, consequently, the icy particles of the cometary halo may fully consist of such solid mixture or, at least, contain a significant amount of it. Solar UV radiation can excite photoluminescence of icy particles in the halo. A low albedo of the particles, consisting of those mixtures and a high quantum yield of photoluminescence for PAHs ensure detectability of corresponding luminescence emissions.

PAHs and acyclic hydrocarbons, as well as other possible components, remain frozen at low temperatures, and can be therefore named Frozen Hydrocarbon Particles (FHPs).

2. FHPs as carriers of unidentified cometary emissions

Quantum yield of luminescence for many organic compounds is about 0.5. d'Hendecourt *et al.* (1986) stated that the quantum yield for PAHs molecule may be about 50%. According to Gudipati *et al.* (2003) the quantum yield varies from 90 to 100% for smaller grains containing frozen organic mixtures.

The above facts are evidence that, when exposed to UV radiation, frozen organic mixtures have highly efficient, bright luminescence. The cometary nuclear ice may contain mixtures of frozen PAHs and acyclic hydrocarbons, and the mixtures are supposed to appear as a solid solution of the "solvent-substance" type, where the PAHs make the "substance" and the acyclic hydrocarbons are the "solvents". It actually means that PAHs dissolved in acyclic hydrocarbons make a solid solution. The optical properties of such a frozen substance, i.e., polycrystalline solution, depend on the properties of the solvent, solution crystallization conditions, presence of a luminescent component, a luminogen, relation between the components of the solution and the composition of admixtures in the solution.

Presence of such polycrystalline solutions in the cometary ice is rather probable. The surface layers of an icy cometary nucleus are the sources of heterodispersive frozen hydrocarbon particles, ejected and carried into the circumnuclear area while the heliocentric distance of the comet diminishes. These icy particles of different sizes and irregular symmetry will form an icy halo at corresponding heliocentric distances. We believe that the size of an individual FHP may vary from microns to millimeters. FHPs will have a specific colour characterizing frozen PAHs+acyclic hydrocarbons mixture. With low concentrations of admixtures, FHPs should appear as gray icy particles. So, cometary nuclei can be surrounded by a halo of an FHP layer. In natural conditions, in addition to FHPs, a halo may contain silicate dust and finely dispersive carbonaceous particles. Exposed to solar UV radiation, FHPs produce intensive luminescence within 3800–6700 Å. Other components of halo, such as inorganic dust particles, will also luminesce, though with low quantum yield. FHP luminescence spectrum depends on the chemical composition of a particular mixture, PAH concentration in the polycrystalline solution, admixtures, if there are any, temperature and, occasionally, size of the icy particles, solar activity phase and some other factors, among which the chemical composition and temperature of a particular FHP are the most important.

When exposed to solar UV-radiation, FHPs of different chemical compositions and temperatures will produce luminescence of different spectral compositions. Pringsheim *et al.* (1951) noted that at 54.3 K (temperature of liquid oxygen), diffuse spectral bands of many crystalline organic compounds resolve into groups of narrow bands or lines. This is particularly significant for FHPs and suggests of, at least, two types of luminescence spectra of the cometary halo icy particles: (a) spectra with wide diffuse bands, and (b) spectra in the form of a series of narrow lines or bands.

Let us now dwell on some properties of FHPs in a 60–85 K temperature range, referring to laboratory data of FHP substance analogs. Teplitskaya *et al.* (1978) obtained polycrystalline solutions of the above described composition to study PAH luminescence at 77 K, and found normal paraffins to be the best solvent in producing PAH

discrete luminescence and absorption spectra. In the case of compounds with a linear structure (polyacene, polyphenyl, diphenylpolyene, etc.) the sharpest spectra were observed with the linear dimensions of the solvent molecules close to those of the PAH molecules.

Experimental excitation of an aromatic hydrocarbon solution with laser light in (0–0) band at $T = 4.2$ K produced high-structure hydrocarbon spectra with the line width of 0.15–0.47 Å. This is true for majority of the solvents that crystallize and glassify while freezing, remaining chemically neutral to the molecules solved therein, and optically transparent in the absorption and emission areas of PAH molecules. Experiments also demonstrated that at low PAH concentrations in some solutions, the mixture luminescence spectrum consists of a band series, while at higher PAH concentrations a series of narrow lines may form. Luminescence spectrum of pyrene at its concentration $10^{-4} \text{ mol} \cdot \ell^{-1}$ in *n*-pentane did not reveal a narrow line. Shpol'skii (1962) proved that when PAHs were dissolved in special solutions – hydrocarbons (e.g., in *n*-pentane) at 77 K and lower temperatures, the resulting frozen mixture produced a luminescence spectrum of a series of numerous very narrow lines. Such spectra are called luminescence quasilinear spectra. Position in the band length scale, relative spacing and relative intensity of the lines are fairly specific for each molecular structure, indicating normal electronic-vibrational state of the corresponding molecular structures. We suppose that FHPs cometary halos which are chemically identical to their laboratory analogs, also produce quasilinear luminescence spectra when exposed to UV radiation.

It is noteworthy, however, that FHP quasilinear luminescence spectra produced in a natural cosmic environment may differ from laboratory results due to specific cometary coma temperature and pressure, bombardment with charged particles, or collisions with gaseous neutrals of cometary atmosphere. The width of each FHP luminescence spectral line seems to depend on the heliocentric distance, since the cometary substance temperature changes with the distance from the sun. Lines of quasilinear luminescence spectra of FHP may demonstrate widening at the substance temperature $T > 85$ K and low PAH concentrations. Gradual widening of the lines in the icy particle luminescence spectra may proceed with a diminishing heliocentric distance, as the FHP volatile component gets exhausted. Thus, if the FHP substance consists of PAHs+*n*-hexane, the depletion process starts as soon as the icy particle's temperature reaches $T \approx 290$ K, at which point the *n*-hexane begins to evaporate vigorously. This will naturally change the luminescence spectrum of respective FHPs. Consequently, FHPs can form as sources of short-term luminescence. Such short-lived emissions are not infrequent in cometary spectra and are often unidentifiable. So, the FHP luminescence spectrum for cometary halo depends on:

- the chemical composition of the frozen mixture,
- PAH concentrations,
- heliocentric distance,
- dynamic properties of the halo.

One should also notice the role of the exciting radiation since its range and intensity are considerably different from laboratory ones. The icy particles in the halo are exposed to a short-wave solar radiation of a fairly wide range, and fluxes of charged particles of the solar wind.

Photoprocessing of FHPs exposed to solar UV radiation affects the FHP luminescence spectrum, changing its composition as well as distribution of the luminescence emission intensities in the series. It is worth-while mentioning that those changes occur during a certain time interval that could be defined as an FHP photoprocessing period, which may vary from a few hundred seconds to a few years for each class of FHPs at different heliocentric distances.

The whole picture appears to be complicated, changeable, dependent on the icy particle temperature, PAH concentration and composition, exciting radiation composition. Nevertheless similar photoluminescence emissions are certain to be recorded from different comets at the same heliocentric distances (e.g., $r = 1$ a.u.) with assumed chemical identity of the halo icy particles. So, spectra of different comets may demonstrate similar emissions that have so far been considered unidentifiable.

We suggest interpreting some of the unidentified emission lines in cometary spectra as the photoluminescence of FHPs. We performed calculations of the ratio of the photoluminescence flux to the scattered solar radiation flux F_{lum}/F_{scat} for typical conditions of actual cometary halos composed of millimeter-size FHPs. The calculation was performed on the following conditions: the FHP is composed of phenanthrene $C_{14}H_{10}+n$ -hexane at $T = 77$ K, and the phosphorescence line is $L = 4602 \text{ \AA}$ (Teplitskaya *et al.* 1978). We considered an FHP that scattered solar radiation with a wavelength of $L = 4602.17 \text{ \AA}$. According to Cochran *et al.* (2002) in the spectrum of comet 122P/de Vico was observed unidentified emission at $\lambda = 4602.17 \text{ \AA}$. Subsequent calculations were performed using Plank's formula and an expression for energy exposure $E = (w/s)/r^2$ (Simonia *et al.* 2003) with $w = F \cdot t$ (flux and time) where w is the total energy emitted at a corresponding wavelength, s is the halo's surface area with radius $R = 500$ km, r is the heliocentric distance (1 a.u.). The assumed FHP luminescence quantum yield was 50%, the FHP albedo $A = 0.1$, the exciting UV solar radiation wavelength was $2930.25 \text{ \AA}$ (phenanthrene absorption area). We found $F_{lum}/F_{scat} = 2.344$, which proves that the luminescence signal lies above the scattered solar continuum and is a weak but quite recordable emission. Such relatively weak unidentified emissions are widely encountered in the atlas by Cochran *et al.* (2002). (The error of $\pm 0.17 \text{ \AA}$ is negligible, particularly in view of the above described widening of the corresponding lines.) A high quantum yield of the PAH luminescence and application of fast instrument and high resolution spectrographs ensure that the luminescent signal is recognizable. We calculated also F_{lum}/F_{scat} ratios for other emissions for the same comet. We obtained similar results $F_{lum}/F_{scat} > 1$.

We compared laboratory luminescence spectra of FHP substance analogs with the observed cometary spectrum containing a set of unidentified emissions. We used an atlas of quasi-line luminescence spectra for aromatic molecules (Teplitskaya *et al.* 1978) as the laboratory data base and a High Resolution Atlas of Comet 122P/de Vico (Cochran & Cochran 2002) as the observation data. The results were summed up in a table, a column of the table lists the PAH descriptions and formulas, b column gives the solvent, c column the luminescence emission wavelengths for polycrystalline solutions and d column cites the unidentified cometary emission wavelengths from de Vico spectrum. The allowed error of the analysis was $\pm 1 \text{ \AA}$.

a	b	c	d
Pyrene C ₁₆ H ₁₀	<i>n</i> -hexane	3929	3929.28
		5894	5893.85
		6027	6026.66
		6036	6036.42
		6346	6346.78
		6413	6413.07
		6509	6508.71
1,2-3,4 dibenzanthracene C ₂₂ H ₁₄	<i>n</i> -octane	3862	3862.07
		3936	3936.23
		3949	3949.15
		4069	4069.03
		4091	4090.99
	<i>n</i> -hexane	3934	3934.90
		3955	3955.13
		3981	3981.64
		4055	4054.64
		4067	4067.35
Anthracene C ₁₄ H ₁₀	<i>n</i> -hexane	3950	3950.02
		3963	3963.11
		3985	3985.44
		4011	4011.44
		4023	4023.07
		4075	4075.98
		4248	4248.06
		4264	4263.07
		4292	4291.73
		4490	4489.90
	<i>n</i> -heptane	4519	4519.09
		3867	3867.16
		3967	3967.51
		3987	3987.26
		4002	4001.97
		4025	4024.93
		4064	4064.46
		4094	4093.94
		4295	4294.88
		4541	4540.82
Diphenyl C ₁₂ H ₁₀	<i>n</i> -hexane	4573	4572.97
		4371	4371.16
Fluorene C ₁₃ H ₁₀	<i>n</i> -hexane	4570	4570.12
		4239	4238.71
		4525	4525.13
Phenanthrene C ₁₄ H ₁₀	<i>n</i> -hexane	4552	4552.25
		4591	4591.28
		4602	4602.17
		4957	4957.43
		4991	4991.97
		5060	5060.23
		5313	5313.29
		5327	5327.89
Chrysene C ₁₈ H ₁₂	<i>n</i> -hexane	5400	5400.20
		4000	4000.09
		4983	4982.63
		5028	5028.42
		5058	5057.71

a	b	c	d
		5318	5318.08
		5352	5352.12
		5422	5422.38
		5510	5509.39
Coronene C ₂₄ H ₁₂	<i>n</i> -hexane	4330	4329.69
		4434	4433.11
		4451	4451.59
		4536	4536.33
		4543	4543.00
		4550	4549.99
		5159	5159.89
		5260	5260.31
		5510	5509.39
		5615	5615.23
		5630	5629.74
		5930	5929.59
Tetracene C ₁₈ H ₁₂	<i>n</i> -hexane	4751	4750.96
		5087	5086.61
		5238	5237.57
	<i>n</i> -nonane	4719	4719.17
		4794	4794.01
		4866	4866.26
		5135	5135.96
		5176	5176.26
		5197	5197.21
		5219	5218.65

Described FHP concept and the comparative analysis of the observation with laboratory data summed up in the above table suggest that an FHP of the ice halo of de Vico comet may comprise, as its basic mixture, few aromatic hydrocarbons producing a polycrystalline solution with a number of acyclic hydrocarbons. It is only natural for an FHP to contain small amounts of admixtures. In fact, de Vico's icy halo can be a set of FHPs of various chemical compositions. A cloud of heterodispersive frozen hydrocarbon particles of different chemical compositions exposed to solar UV radiation produces luminescence emissions within a wide range of 3800–6700 Å. Our comparisons involved laboratory data on hydrocarbon fluorescence and phosphorescence. Polycrystalline solutions – chemical analogs of the FHP substance demonstrated in the laboratory intensive fluorescence and phosphorescence. In the natural cosmic environment with different excitation conditions, FHP of the cometary halo may have different afterglow periods, i.e., they may produce fluorescence or phosphorescence spectra. The laboratory base that was used in our comparative study contained 100 aromatic molecules. Occurrence of PAHs in comets was a subject of study for quite a few authors, for instance, Moreels *et al.* (1994) proved that the inner coma of comet P/Halley contains phenanthrene. They detected this aromatic hydrocarbon in the near UV range of the comet spectrum. They pointed out stability of PAHs in the cosmic environment and suggested possible relation between PAHs of the solar system bodies and PAHs of interstellar origin. According to Crovisier *et al.* (1999) the presence of PAHs in comets still remains doubtful. Ehrenfreund *et al.* (2000) mentioned detection of PAHs and aliphatic hydrocarbons in the cometary substance in the section “Cometary Dust and Refractory Organics” of their extensive review.

Aromatic hydrocarbons can be found in comets both in the condensed and gaseous phase. The above described FHPs of a corresponding chemical composition carry a number of photoluminescence emissions that have been so far considered as unidentified ones, although it is hardly possible that all the unidentified emissions are actually FHP photoluminescence. Other parts of unidentified emissions belong to resonance fluorescence of daughter molecules and ions. The Flum/Fscat ratio was calculated assuming the albedo value of 0.1.

3. Conclusion

The concept of FHPs has been suggested and basic features of their photoluminescence have been described. It was proposed to regard at least a part of the unidentified cometary emissions as FHP photoluminescence. A comparative analysis was carried out involving spectra of FHP's substance laboratory analogues with spectra of de Vico comet containing several thousands of unidentified cometary emissions, which demonstrated that the icy particles of that comet's halo may contain at least 9 aromatic molecules. This work is only the first step in this field of research, and our intention is to develop it further by updating our methods and results.

References

- Brown, M. E., Bouchez, A. H., Spinrad, H., Johns-Krull, C. M. 1996, *Astron. J.*, **112**, 1197.
Cochran, A., Cochran, W. 2002, *Icarus*, **157**(2), 297.
Crovisier, J., Bockelee-Morvan, D. 1999, *Space Science Reviews*, **90**, 19.
Ehrenfreund, P., Charnley, S. B. 2000, *Annu. Rev. Astron. Astrophys.*, **38**, 427.
d'Hendecourt, L. B., Leger, A., Olofsson, G., Schmidt, W. 1986, *Astron. Astrophys.*, **170**, 91.
Gudipati, M. S., Dworkin, J. P., Chiller, X. D. F., Allamandola, L. 2003, *Astrophys. J.*, **583**, 514.
Moreels, G., Clairemidi, J., Hermine, P., Brechignac, P., Ronsselot, P. 1994, *A&A*, **282**, 643.
Pringsheim, P. 1951, *Fluorescence and Phosphorescence*, (Moscow).
Shpol'skii, E. V. 1962, *UFN*, **77**.
Simonia, I. 1999, *Comments on Modern Physics, Comments on Astrophysics*, **1**(E), 25.
Simonia, I., Simonia, T. 2003, *Astron. Astrophys. Trans.*, **22**(1), 55.
Teplitskaya, T. A., Alexeeva, T. A., Valdman, M. M. 1978, *An Atlas of Quasilinear Luminescence Spectra from Aromatic Molecules*, (Moscow University).

Average Anisotropy Characteristics of High Energy Cosmic Ray Particles and Geomagnetic Disturbance Index Ap

C. M. Tiwari¹, D. P. Tiwari¹, Ajay K. Pandey¹ & Pankaj K. Shrivastava²

¹*Department of Physics, A.P.S. University, Rewa (MP) 486 003, India.*

e-mail: tiwaricm123@rediffmail.com

²*Department of Physics, Govt. Model Science College, Rewa (MP) 486 001, India.*

e-mail: pankaj_in_2001@rediffmail.com

Received 2003 October 15; accepted 2005 April 28

Abstract. The average characteristics of the diurnal and semi-diurnal anisotropy of cosmic ray intensity at relativistic energies have been obtained by using data from the worldwide grid of neutron monitor for the period 1989 to 1996. The complex behaviour of the diurnal amplitudes and time of maxima (phase) and its association with the Ap index on a long-term and day-to-day basis have been studied. Even though the general characteristics, on a yearly average basis, have not changed significantly during this period, both the diurnal and semi-diurnal amplitudes and phases vary significantly, besides significant changes being observed for different interplanetary conditions on a short-term basis. It is found that the relationship between the Ap index and the diurnal vector is out of phase during the period 1991 to 1995. On a long-term basis, the correlation of diurnal variation with Ap index has been found to vary during the solar cycle. On a short-term basis, it has been observed that the high Ap days are usually associated with higher amplitudes with phase shifted to earlier hours.

Key words. Cosmic rays—anisotropic variation—geomagnetic disturbances.

1. Introduction

Long-term variability (11 yr/22 yr) periodicity in diurnal variation of cosmic ray intensity was proposed as early as in 1953 (Thambyahpillai & Elliot 1953). Changes in amplitudes and time of maxima (phase) have been studied by a number of investigators (Rao 1972; Forbush 1973; Agrawal 1983; Shrivastava 1990; Tiwari 2001; Tiwari *et al.* 2003). Characteristics of daily variation were also investigated on a day-to-day basis for different interplanetary conditions using Ap index for the interval period 1968 to 1974 (Sharma *et al.* 1978). Ap value is a geomagnetic disturbance index, which is recorded by magnetometers/magnetograms. The quantitative values of the geomagnetic field are obtained by formulating the geomagnetic indices. A characteristic of the study of geomagnetic field variation of magnetic activity over the earth surface is a complicated task. However, this has been done through the use of magnetic indices (Mayaund 1980). Ap and Kp are examples of such indices, which provide physical

significance of field variation. Shrivastava & Shukla (1993) have reported a significant influence of high-speed solar wind streams on geomagnetic activity enhancement. Further Shrivastava & Shukla (1996) reported that there is a high correlation between solar wind velocity and Ap index. As we know from convection diffusion approximate theory, solar wind velocity plays an important role in cosmic ray modulation. In the absence of solar wind data, one can use the daily values of Ap index. Since Ap values are well correlated with solar wind velocity, we can use Ap values to explain the anisotropic variation in border areas. High- and low-Ap values also indicate disturbed and quiet interplanetary medium. In this paper, we have critically examined the relationship of diurnal/semi-diurnal amplitude and phases with geomagnetic disturbance index (Ap) for the period 1989 to 1996.

2. Results and discussion

The temperature and pressure corrected hourly data (counts of neutrons) of cosmic ray intensity from two neutron monitors have been used to obtain the amplitude and phase of the harmonic component for each day by Fourier techniques, where the long-term change from the data has been removed by the method of 24-hour moving average. The days of Forbush decreases have also been removed from the analysis to avoid associated superposed variation in cosmic rays. A complete computer program is made to calculate the harmonics of every day for the period 1989 to 1996. The yearly mean values with their standard errors have been obtained from these daily vectors. Geomagnetic disturbance index Ap is found to be a very good proxy of the interplanetary conditions. The low values of Ap index of geomagnetic activity indicate quiet periods, thereby signifying quiet conditions of the interplanetary medium. For this we have chosen (1) days for which $Ap \leq 8$ and derived the amplitude and phases for such days, which are again averaged to get the yearly mean value. Likewise other criteria are Ap values ranging between 9–17, 18–26, 27–53 and $Ap \geq 54$. Higher Ap groups are an indicator of disturbed interplanetary medium.

As mentioned earlier the statistical relationship between Ap and amplitude (r_1) and also with time of maximum (ϕ_1) of diurnal variation of all on a day-to-day basis was found to change drastically, during the period 1968 to 1973 (Sharma *et al.* 1978; Agrawal *et al.* 1995). In fact, it has been demonstrated that the Ap index is highly correlated with the mean fluctuations in the amplitude of the interplanetary magnetic field, which in turn is related to the diffusive component of the convection–diffusion theory. Moreover, Ap is also found to be related to the solar wind velocity, which is related to convective component (Sabbah *et al.* 1996). It is therefore, natural to look for the relationship between Ap and diurnal/semi-diurnal variation for the recent period also.

To obtain the relationship between Ap index with diurnal and semi-diurnal anisotropy on a day-to-day basis, we have taken the cosmic ray data from two stations Kiel (cut-off rigidity ≈ 2.32 GV) and Tokyo (cut-off rigidity ≈ 11.50 GV) covering the high and low cut-off rigidity. The vector average of diurnal and semi-diurnal variation has been obtained by dividing all good days in five groups of Ap values. To understand the variational effects of the solar output during the period of study, the analysis has been extended on a day-to-day basis by using the daily geomagnetic Ap index as the proxy for interplanetary disturbances. For each year, the averages for the diurnal variation for five groups of Ap days have been calculated, and are shown in Fig. 1, for each year from 1989 to 1996. Hourly count rates of cosmic ray neutrons

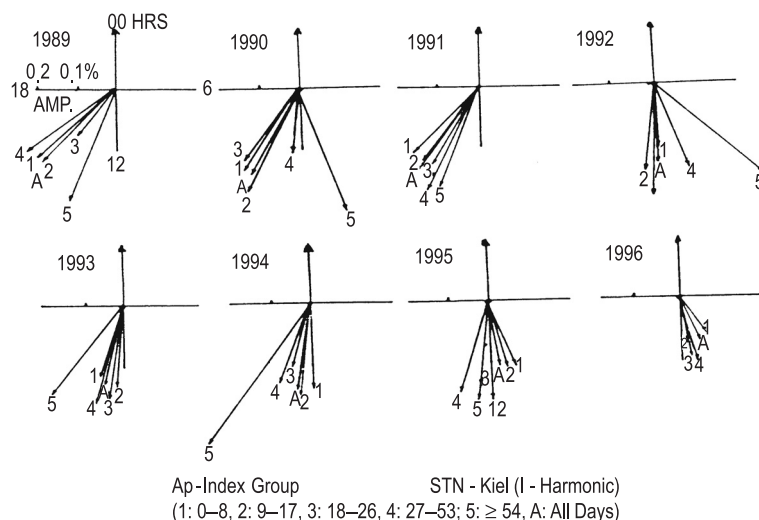


Figure 1. Shows the harmonic dial representation of the diurnal vectors (first-harmonic) for various groups of Ap days, for each year from 1989 to 1996. Group 1 of Ap days comprise of most quiet days, whereas Ap days in group 5 are mostly disturbed days. The average diurnal vectors for each year for all “good days” represented by “A” are for Kiel neutron monitoring station.

are monitored at Kiel and Tokyo neutron monitor stations. These 24-hour counts are used to derive the amplitude and time of maxima by adopting the harmonic analysis techniques. The divisions of days are according to increasing values of Ap index (0–8, 9–17, 18–26, 27–53, and $Ap \geq 54$). However, there is no significance of the fifth group as these days are highly disturbed days. Figure 1 shows the change in diurnal variation with Ap for Kiel station and demonstrates exactly similar variation except the years 1992 and 1996. Further it reveals a complete reversal of the vector from 1 to 4 as the period of maxima to minima approaches. It is seen that in the years 1989 to 1992 and with the increasing values of Ap index the time of maximum shifts to earlier hours, with amplitudes remaining almost constant (except for the group of most disturbed days – group 5), whereas during 1993 to 1995 the situation is just reverse. Such behaviour was also reported during the interval 1968 to 1974. It was reported that in the year 1969 (high sunspot activity period), with the increasing value of Ap index, the time of maximum shifts to earlier hours, whereas during 1973, it was found to shift to later hours (Agrawal & Singh 1975). This indicates that once again the interplanetary conditions must be drastically changing during this period of our study. Figure 2 shows the result of similar analysis for Tokyo stations. The vector average values for each group and for each year from 1989 to 1996 for Tokyo stations are plotted in this figure. It is observed that for low Ap group the dispersion from one period to another is least in both the amplitude and phase. However, the dispersion is large for two higher Ap index groups particularly for Ap, 27–53 and $Ap \geq 54$ (groups 4 and 5). It is also noticed that the amplitude (r) increases with Ap result consistent with earlier findings (Sharma *et al.* 1978). Nevertheless, it appears that the vectors associated with high Ap values having large amplitude and large scatter-diagrams are composed of real extra-terrestrial anisotropy and that the variation of terrestrial origin which could be due to universal time variations, cut-off rigidity changes correlated

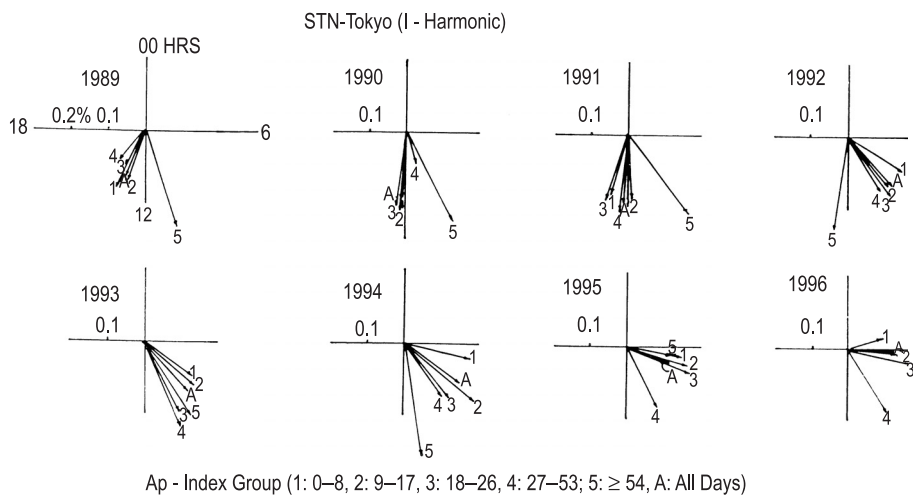


Figure 2. Similar to Fig. 1, but for Tokyo neutron monitor data.

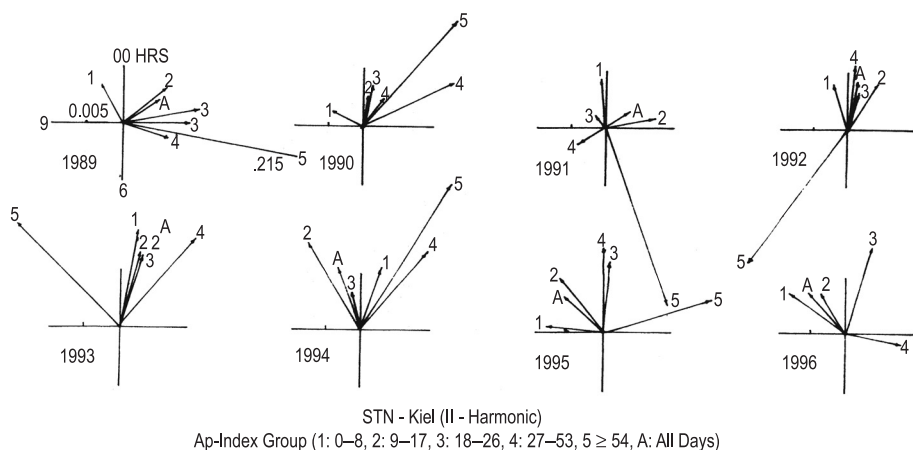


Figure 3. Shows the harmonic dial representation of the semi-diurnal vectors (second-harmonic) for various groups of Ap days, for each year from 1989 to 1996. Group 1 of Ap days comprises of most quiet days, whereas Ap days in group 5 are mostly disturbed days. The average diurnal vectors for each year for all “good days” represented by “A” are for Kiel neutron monitoring station.

with geomagnetic field variations, and/or due to inadequate atmospheric pressure correction. During the year 1989–1992, the diurnal time of maximum is found to shift to earlier hours with the increasing values of Ap index, which is in complete agreement with earlier findings (Agrawal & Singh 1975). The complete reversal of the nature of Ap and ϕ_1 during the years 1989–1992 and earlier, is not very suppressing due to the fact that Ap has increased during 1989–1992 though it was expected to decrease in this period, and that ϕ_1 has shown a large change is not seen. It may be noted that the relationship is not so clear during most of the years. Figure 3 shows the similar harmonic dial for semi-diurnal vectors in all five groups for Kiel stations cosmic ray data. Similarly, Fig. 4 shows the results for Tokyo neutrons. Large variations in

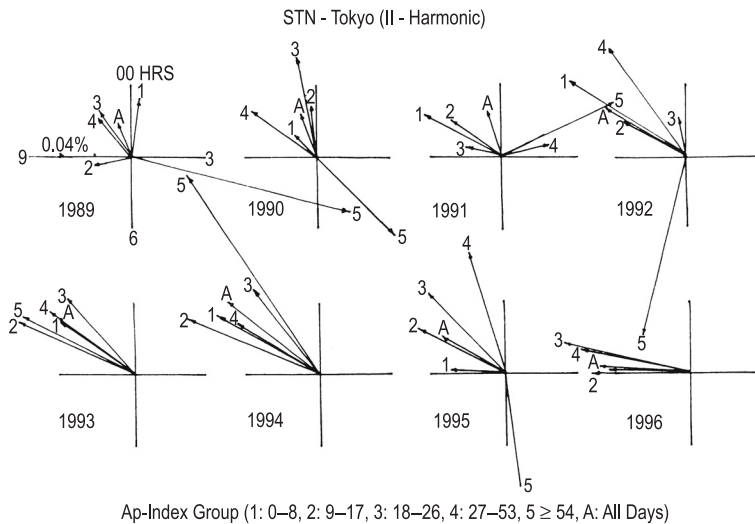


Figure 4. Similar to Fig. 3, but for Tokyo neutron monitor data.

average vectors for all the five groups are seen. Moreover, no definite trend in the semi-diurnal variation with Ap is established. However, increase in its amplitude is found with the increasing value of Ap index. The results of the analysis reported here has two components: firstly the long-term variation of the diurnal anisotropy, i.e., the change in the overall values of the annual average diurnal amplitude and phase. It is observed that in a typical high/middle latitude neutron monitor station such as Kiel, the diurnal amplitude is substantially high during the maximum phase of the sunspot cycle, whereas the diurnal amplitude is quite small during decreasing phase and more so in the minimum phase of the sunspot cycle. Secondly, when the days are distributed as per the degree of interplanetary disturbances (proxy based on Ap), one finds that the behaviour of the diurnal variation is quite different during the two periods mentioned above. The diurnal time of maximum rotates to earlier hours during maximum phase whereas it rotates to later hours during the declining phase of the solar activity. The two contradictory situations can be qualitatively understood in terms of two different types of interplanetary conditions prevailing during different phases of the sunspot activity cycle.

It is now well known that there are two types of solar wind streams, one associated with solar flares and the other associated with solar coronal holes occurring in exclusion with each other, i.e., the first occurring abundantly during maximum phase of sunspot activity, whereas the other, coronal holes associated streams, abundantly occurring during declining phase of the solar activity. Such a difference can be associated with the peculiar observations reported here, however, the mechanism is not yet fully understood.

In this context, it is also pertinent to note that the short-term cosmic ray intensity variations have been reported to be quite different for the two types of solar wind streams. The coronal hole associated high speed solar wind streams do not produce any large decrease in cosmic ray intensity, in contrast to very large Forbush type cosmic ray decreases associated with high velocity solar flare streams (Venkatesan *et al.* 1982; Shrivastava & Shukla 1994, 1996). The spectral exponent of

the cosmic ray variation was also reported to be quite different in the two cases. As such the nature of observed anomalous changes in the diurnal variation and their relation with Ap index does not reveal the exact mechanism, which is operative during the period of study.

The results may lead to an easy way out to predict the occurrence of the solar phenomena from ground-based cosmic ray observations.

3. Conclusions

The analysis presented leads to the following conclusions:

- It is found that in diurnal anisotropy for low Ap group of dispersion from one year to another is least particularly in amplitude (Ap values ranging between 27 to 53 and $Ap \geq 54$). However, the dispersion is large for high Ap values (groups 4 and 5).
- Time of maxima (phase) for diurnal anisotropy shifts to earlier hours for the period 1992 to 1996.
- The semi-diurnal amplitude and phase have not changed very significantly except in later periods associated with high values of Ap index (groups 4 and 5).
- The semi-diurnal amplitude and phase associated with low Ap values are more coherent and are probably representative of the true extra-terrestrial anisotropy, whereas large increase in amplitudes and variability during high Ap days could be related to terrestrial origin.

Acknowledgements

Authors are thankful to Dr. Alexai Struminski of Moscow University, Russia for his suggestions and for providing the hourly neutron monitor data. Authors are also thankful to Prof. S. P. Agrawal, Ex Vice-chancellor, A.P.S. University, Rewa (MP) for helpful discussion and suggestions.

References

- Agrawal, S. P., Shrivastava, P. K., Shukla, R. P. 1995, *Proc. 24th Int. Cos. Ray Conf. Italy*, **4**, 932.
 Agrawal, S. P. 1983, *Space Sci. Rev.*, **34**, 127.
 Agrawal, S. P., Singh, R. L. 1975, *Proc. 14th Int. Cos. Ray Conf. Munich*, **3**, 1253.
 Forbush, S. E. 1973, *J. Geophys. Res.*, **78**, 7933.
 Forman, M. A., Gleeson, L. J. 1975, *Astrophys. & Space Sci.*, **32**, 771.
 Mayaund, P. N. 1980, Geographical Monograph, 22 AGU, Washington, D.C., USA.
 Rao, U. R. 1972, *Space Sci. Rev.*, **12**, 719.
 Sabbah, I., El-Borie, M. A., Pandey, A., Shrivastava, P. K. 1996, *Indian J. Phys.*, **70B(5)**, 367.
 Sharma, R. M., Nigam, S. K., Singh, R. L., Agrawal, S. P. 1978, *Indian J. Radio & Space Phys.*, **7**, 210.
 Shrivastava, P. K., Shukla, R. P. 1996, *Bull. Astr. Soc. India*, **24**, 663.
 Shrivastava, P. K. 1990, *Proc. 21th Int. Cos. Ray Conf. Adelaide*, **6**, 353.
 Shrivastava, P. K., Shukla, R. P. 1993, *Proc. 23rd Int. Cos. Ray Conf. Calgary, Canada*, **3**, 590.
 Shrivastava, P. K., Shukla, R. P. 1994, *Sol. Phys.*, **154**, 177.
 Shrivastava, P. K., Shukla, R. P. 1996, *J. Astron. Astrophys.*, **16**, 297.
 Tiwari, C. M. 2001, Ph.D. Thesis, A.P.S. University, Rewa, India.
 Tiwari, C. M., Agrawal, S. P., Tiwari, D. P., Elbori, M. A., Shrivastava, P. K. 2003, *Curr. Sci.*, **3(1)**, 219.
 Thambyab Pillai, T., Elliot, H. 1953, *Nature*, **171**, 918.
 Venkatesan, D., Shukla, A. K., Agrawal, S. P. 1982, *Sol. Phys.*, **81**, 375.

Small-scale Magnetic Field Diagnostics outside Sunspots: Comparison of Different Methods

D. N. Rachkovsky¹, T. T. Tsap¹ & V. G. Lozitsky²

¹*Crimean Astrophysical Observatory, Nauchny, Crimea, UA-98409, Ukraine.*

²*Kyiv University Astronomical Observatory, Observatorna St. 3, Kyiv, UA-04053, Ukraine.
e-mail: lozitsky@observ.univ.kiev.au*

Received 2004 December 28; accepted 2005 November 28

Abstract. We analyse different observational data related to the problem of intrinsic magnetic field strength in small-scale fluxtubes outside sunspots. We conclude that the kG range of fluxtube fields follows from not only classical line ratio method, but also from other old and new techniques. For the quiet regions on the Sun, the most probable mode of such fields has a magnetic field strength of 1.2–1.5 kG assuming the rectangular field profile. To best interpret the observations, a weak background field between fluxtubes should be assumed, and its magnetic field strength is expected to increase with the filling factor of fluxtubes. The alternative point of view about subkilogauss fluxtube fields is critically examined, and possible sources of different conclusions are presented.

Key words. Sun—magnetic fluxtubes—kG magnetic fields.

1. Introduction

Recently Zirin & Cameron (2002) published a very interesting paper with new observational data concerning the problem of intrinsic magnetic field strength of non-spot regions. These authors proposed a new method of magnetic field diagnostics for regions of weak measured fields which is similar to the line ratio technique proposed by Stenflo (1973). The main conclusion of Zirin & Cameron (2002) is that ‘the invisible kilogauss fields postulated by Stenflo do not exist in these weak field elements’. Note that this conclusion contradicts not only J. O. Stenflo, but also many other investigators in the field of magnetography. So, we have at hand, a very interesting problem, which would influence our understanding of the physical nature of the small-scale magnetic fields on the Sun.

Detailed magnetic characteristics of the small-scale magnetic elements are practically unknown due to essential problems with the magnetic field measurements in very inhomogeneous solar plasma. This problem has been discussed by many authors (Severny 1964, 1965, 1966, 1967; Steshenko 1967; Howard & Stenflo 1972; Stenflo 1973; Frazier & Stenflo 1978; Wiehr 1978; Koutchmy & Stellmacher 1978; Lozitsky 1980; Rachkovsky & Tsap 1985; Staude & Hoffman 1988; Carroll & Staude 2001, etc.).

Firstly, the most important observational clue on the existence of the spatially unresolved magnetic fields with kG fields was discovered with a Babcock (1953) type

solar magnetograph, in two lines of the same multiplet of FeI, 5247.1 and 5250.2 Å, which have practically equal levels of formation and temperature sensitivity, but different Lande factors g , 2.0 and 3.0, respectively (Stenflo 1973). Theoretically, in case of subkilogauss magnetic fields these lines must give practically the same measured field values, as manifested by the longitudinal magnetic field $B_{\parallel}$ in the solar atmosphere. In practice however, the first line gives, as a rule, about 15–20% more strong field than the second one. In addition, the line ratio $B_{\parallel}(5247)/B_{\parallel}(5250)$ was observed to increase from line center to its wing. On the basis of calculations, Stenflo (1973) showed that this anomalous behaviour of the line ratio indicates the intrinsic magnetic field strength to be in the range 1.1–2.3 kG. Although he observed with spatial resolution ~ 1700 km in quiet regions on the Sun, he postulated the existence of the invisible magnetic flux tubes with probable diameters 100–300 km.

This was a very courageous idea, which was proposed in the same time when the problems of magnetographic signal calibration were not fully resolved. In particular, in the same year Gopasyuk *et al.* (1973) published interesting results about magnetic field comparison in different spectral lines. It was shown for about ten lines, that, in general, measured field $B_{\parallel}$ decreased with the magnetic sensitivity factor $g\lambda^2$. As pointed out later by Lozitsky (1979, 1980), the observed peculiarities of this dependence can not be explained in terms of spatially unresolved magnetic fields with intrinsic strengths of 1–2 kG. An alternative suggestion was that this dependence had an instrumental origin.

In addition, Severny (1967) obtained the essential differences between theoretical and experimental calibration curves of the solar magnetograph, using the sunspots as objects on the Sun with well detectable magnetic field. Further, similar unexplained differences were obtained for laboratory calibration of the magnetograph, for the case of a known homogeneous field. This raises the question: if one cannot measure the *homogeneous* fields (in the laboratory) how can we measure the *inhomogeneous* fields (on the Sun)?

For a way out of the situation, the data of full Stokes profiles were needed – similar to Stokesmeter (Harvey *et al.* 1972), or spectral data obtained with a polarization analyser (Kouchmy & Stellmacher 1978; Lozitsky 1980), or Fourier transform spectrometer (Stenflo *et al.* 1984). They confirmed, in particular, that line ratio differences are not instrumental effects. Practically the same dependence of $B_{\parallel}$ versus $g\lambda^2$, as obtained by Gopasyuk *et al.* (1973), was found by Lozitsky (1979, 1980) using the spectral data based on Stokes I profiles of about ten lines. Lozitska *et al.* (1982) showed that empirical calibration curves of the solar magnetograph obtained by Severny (1967) may be explained by more narrow spectral lines of HgI in laboratory spectra, and also by the increase in half-widths that occurs due to increasing of the total pressure (gas + magnetic) in a glass balloon. So, some of named problems were partly resolved, and it supports indirectly the main conclusion about the reality of the strong unresolved magnetic fields on the Sun. In addition, a very wide analysis of the statistical properties of the Stokes I and V line profiles of 400 unblended FeI lines showed that the field strength in network and plages is in the range 1.26–1.72 kG, and filling factor is from 3% to about 15% (Solanki & Stenflo 1984).

In this paper, we analyse the recent results concerned with this problem and consider the most reliable and direct data related to intrinsic magnetic field range in non-spot and non-flare regions.

2. Line ratio method

The basic idea of all methods for unresolved magnetic field diagnostics is the following: if one cannot resolve the superfine structures spatially one could ‘resolve’ spectrally, using an instrument with high spectral resolution and several spectral lines with different Lande factors preferably large in value. In case of an inhomogeneous field (like any fluxtubes with strong magnetic field and weak background field) we should observe the partial spectral polarized contributions in the total observed profile of the line. The position on the line profile of the contributions of different field strengths would depend on the Zeeman splitting rule.

This approach needs the Stokes profiles. On the other hand, the observational data for one point in line wing (like classical magnetograph) or integrated data for wide spectral interval (from line center to continuum – as in filter magnetograph) are practically useless for this goal. This explains why many authors had used several narrow positions of the exit slits of magnetograph in line profiles (Stenflo 1973; Frazier & Stenflo 1975; Wiehr 1978; Rachkovsky & Tsap 1985, etc.).

In the present work we continue to study the small-scale structure of the non-spots magnetic fields using new observational data and new theoretical calculations for these observations. Below we calculate the theoretical diagnostic dependences more exactly than in earlier published papers by Rachkovsky & Tsap (1985), and also Lozitsky & Tsap (1989).

We present here the results of observations made in good seeing near disk center on 24 April 1991 with double Crimean magnetograph (Kotov *et al.* 1982), in FeI 5247.1 and 5250.2 lines. The spectral resolution was 1×2 arc sec, scanned area – 160×160 arc sec related to a quiet region. Three sizes of the exit slits were used: 15–40, 35–60 and 55–80 mÅ from the line center. Magnetograph signals were calibrated and compared in both lines similar to procedure described earlier in detail by Rachkovsky & Tsap (1985).

Theoretical line ratio dependences were calculated using the solutions for the radiative transfer equations in magnetoactive medium (Rachkovsky 1963)

$$\cos \theta \frac{I_{\pm}(\tau, \theta)}{d\tau} = (1 + \eta_0 \sigma_{\pm}) I_{\pm} - (1 - \varepsilon) \eta_0 \frac{3}{8\pi(k_l + k_r + k_p)} \int_{4\pi} (\sigma_+ I_+ + \sigma_- I_-) d\omega - \frac{1}{2} (1 - \varepsilon \eta_0 \sigma_{\pm}) (1 + \beta \tau) B_0, \quad (1)$$

where η_0 is the ratio of line-to-continuum absorption at line center, $\sigma_{\pm}$ are absorption coefficients in spectral line for mutually orthogonal polarizations ‘ $\pm$ ’, ε is probability of full quantum absorption, k_l , k_r , k_p – absorption coefficients for left-, right-, and linear-polarized radiation, β is coefficient of the linear approximation for source function versus optical depth τ in spectral continuum, and B_0 is the Planck’s function.

Mathematical solution for equation (1) depends on the parameter α of Voigt function, i.e., the ratio of damping to Doppler width $\Delta\lambda_D$. These parameters were determined using the observed Stokes I profiles of lines. So long as observed profiles for FeI 5250.2 and FeI 5247.1 lines were found practically the same, line FeI 5250.2 only was used as experimental sample to the theoretical fit (Fig. 1).

One can see from Fig. 1 that theoretical profile (dashed) satisfy the observations (solid line) very well. The following parameters were used: $\eta_0 = 10$, $\varepsilon = 0.05$, $\beta = 2$,

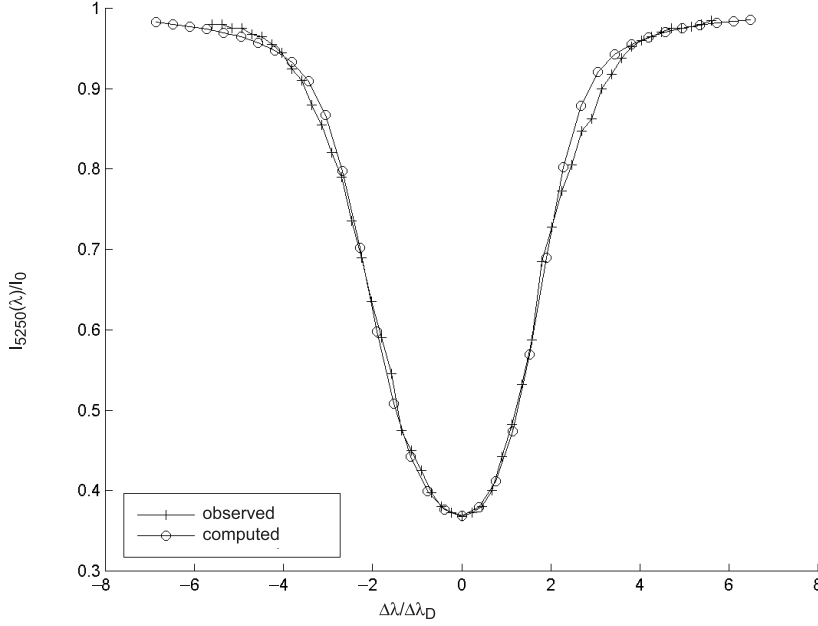


Figure 1. Comparison of the observed and computed profiles for FeI 5250.2 line.

$\alpha = 0.35$, $\Delta\lambda_D = 25 \text{ mÅ}$. The parameter ν is the wavelength displacement from the line center in units of the Doppler width, i.e., $\nu = \Delta\lambda/\Delta\lambda_D$.

Similar to our previous work (Rachkovsky & Tsap 1985), calculations were made for two-component model, according to which the small-scale fluxtubes with strong field B_f are embedded in a more weak background field B_b . Also, we assumed that magnetic field strength was constant in the height range of the line formation in the solar atmosphere, and within the cross-section of each fluxtube (rectangular field profile). In this case,

$$\begin{aligned}
 k &= \frac{S_b \times V_1(B_b) + S_f \times V_1(B_f)}{S_b \times V_2(B_b) + S_f \times V_2(B_f)} \\
 &= \frac{1 + x(V_1(B_f)/V_1(B_b))}{1 + x(V_2(B_f)/V_2(B_b))}, \quad (2)
 \end{aligned}$$

where k is the magnetographic signal ratio for FeI 5250 and FeI 5247, V_1 and V_2 are Stokes V parameters for these lines, respectively, S_b and S_f are relative areas of the weak and strong field, and $x = S_b/S_f$.

The results are given in Fig. 2 and Table 1. The data of the old observations for 1978 are presented as well. We can see all observational data including Stenflo's (1973) measurements are in good agreement.

As it follows from comparison of the observations and theory, all presented observational data need the magnetic fields of *kilogauss range*, namely $B_f = 1100\text{--}1500 \text{ G}$. The probable diameter of the fluxtube was found close to 120 km. As to background field, our calculations do not allow to determine its true strength and possible structure. On the other hand, it is important to note that obtained data strongly indicate

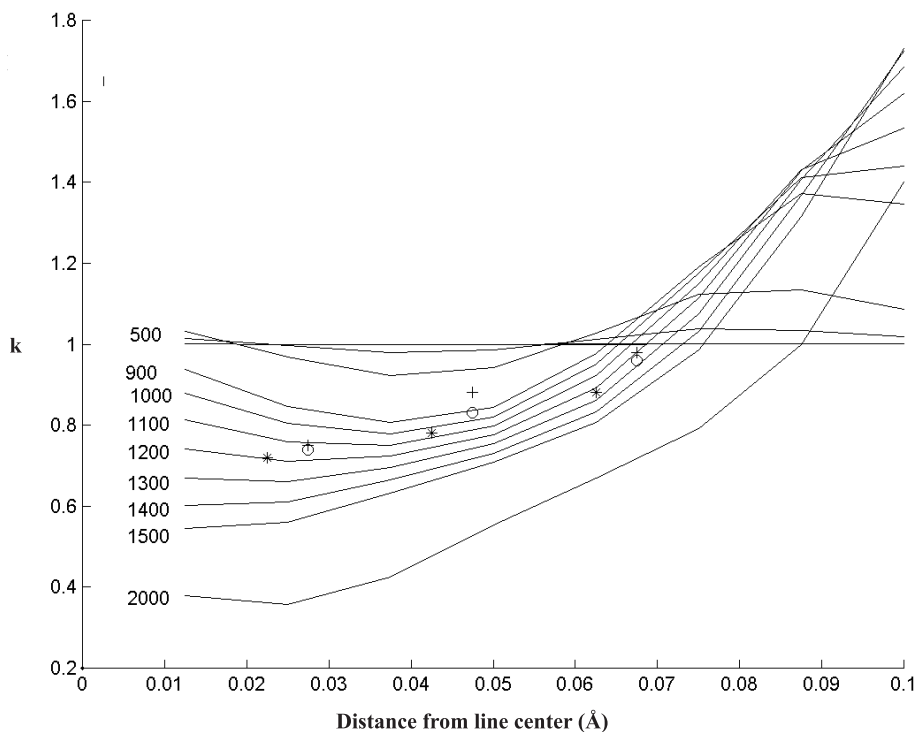


Figure 2. Observed and theoretical dependences for k parameter vs. distance from line center: + – observational data by Stenflo (1973), * and ° – observational data of the present work obtained in 1978 and 1991, respectively, solid curves – theoretical dependences calculated for $\Delta\lambda_D = 22 \text{ mÅ}$, $\eta_0 = 10$, $\varepsilon = 0.05$, $\beta = 2$, $a = 0.35$ and different magnetic field strengths in the magnetic fluxtubes indicated in the figure.

Table 1. Comparison of the observations and calculations for different parameters of the theoretical model.

Exit slit sizes, mÅ	Observations of 1978	Observations of 1991	Model:	Model:	Model:
			$\Delta\lambda_D = 25 \text{ mÅ}$, $x = \infty$, $B_f = 1200 \text{ G}$	$\Delta\lambda_D = 22 \text{ mÅ}$, $x = 0.012$, $B_f = 1300 \text{ G}$	$\Delta\lambda_D = 22 \text{ mÅ}$, $x = 0.006$, $B_f = 1500 \text{ G}$
15–40	0.74	0.75	0.74	0.75	0.74
35–60	0.83	0.88	0.78	0.82	0.81
55–80	0.96	0.98	0.95	0.97	0.92

the existence of the background field. Earlier, it was found that B_f should increase with the filling factor α of fluxtubes, and $B_b/\alpha \approx 1 \text{ kG}$ (Lozitsky & Dolgoplov 1983; Lozitsky & Tsap 1989).

Also, calculations show that for $x = \infty$ the best fit with observations correspond to $\Delta\lambda_D = 25 \text{ mÅ}$, that gives the effective temperature $T \approx 6000 \pm 600 \text{ K}$. One cannot obtain a good agreement between observations and theory for other values of $\Delta\lambda_D$.

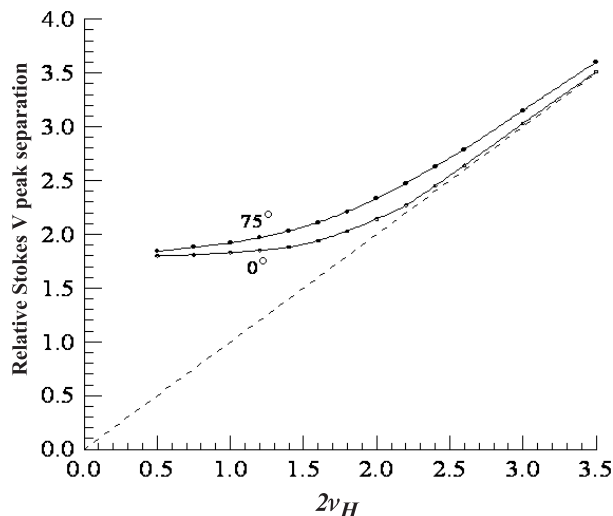


Figure 3. Theoretical dependences of the relative splitting $\Delta\lambda_V/\Delta\lambda_D$ of the Stokes V peaks vs. double relative Zeeman splitting $2\nu_H = 2\Delta\lambda_H/\Delta\lambda_D$ for two angles γ (0° and 75°) between the field line and the line of sight (Lozitska & Lozitsky 1994). It is assumed that $\eta_0 = 0.5$.

So, we can conclude, that line ratio method indicates strongly to the reality of the small-scale magnetic fluxtubes with kG strengths. On the other hand, new observations and theoretical calculations are needed to determine their more reliable characteristics using new methods with more direct Stokes profile recording, and higher spatial and temporal resolution.

3. Stokes V peak separation

The particular drawback of FeI 5250 and FeI 5247 lines is their relatively high temperature sensitivity (Harvey and Livingston 1969; Staude 1970a, b). Lozitska & Lozitsky (1994) have shown that the temperature sensitivity of the pair 5250.2/5247.1 is not an obstacle for reliable magnetic field measurements if we do not use the Stokes V amplitude (such as in a solar magnetograph), but the Stokes V peak separation, $\Delta\lambda_V$. Also, the Stokes V peak half-width, $\Delta\lambda_{1/2,V}$, and the Stokes I half-width, $\Delta\lambda_{1/2,I}$, could be successfully used for small-scale magnetic field diagnostics.

According to calculations, the Stokes V peak separation, $\Delta\lambda_V$, is a very sensitive parameter for Zeeman splitting measurements, especially in case $\Delta\lambda_V/\Delta\lambda_D \geq 2$, or $\Delta\lambda_H \geq \Delta\lambda_D$, where $\Delta\lambda_H$ is the Zeeman splitting (Fig. 3).

As it follows from Fig. 3, for $0.5 < \nu_H < 1.5$ the ratio $\Delta\lambda_V/\Delta\lambda_D$ is a nonlinear function of ν_H . In this interval the field strength cannot be determined directly from the splitting $\Delta\lambda_V$, but from a comparative study of the splitting of lines with different Lande factors.

Corresponding determinations of the value of B is difficult, because $\Delta\lambda_V/\Delta\lambda_D$ depends not only on the parameter ν_H , but also on γ . The latter dependence becomes stronger when η_0 increases. Since there are actually two unknown quantities (B and γ), additional data should be used. Such data can be the half-widths of the V -peaks,

$\Delta\lambda_{1/2,V}$, which also depend on ν_H . This parameter also depends on both ν_H (i.e., B) and γ .

The observations can be interpreted assuming $\nu_H(5250) = 1.5\nu_H(5247)$, while $H(5250) = H(5247)$ and $\gamma(5250) = \gamma(5247)$. This is the idea (firstly proposed by Lozitska & Lozitsky (1994)) behind the line-ratio technique.

On the base of such study of the Stokes V peak peculiarities, the small-scale structure of the solar flares was investigated (Lozitsky *et al.* 1999, 2000). Below we use this method to the non-flare magnetic field diagnostics.

First observational results suitable to this method were obtained by Stenflo *et al.* (1987). In a faculae at the center of the disk $\Delta\lambda_{1/2,I} = 87 \text{ mÅ}$ for FeI 5247 line and 88 mÅ for FeI 5250. On the other hand, for the Stokes V separation, $\Delta\lambda_V$, more essential differences were found, namely 81 mÅ for first line and 100 mÅ for the second. Notice if the magnetic field is really weak and the lines are formed in the undisturbed atmosphere, then we should find $\Delta\lambda_V = 72 \text{ mÅ}$ and $\Delta\lambda_{1/2,I} = 84 \text{ mÅ}$ for both lines (Lozitska & Lozitsky 1994).

Above pointed relations for the faculae $\Delta\lambda_{1/2,I}(5250) \approx \Delta\lambda_{1/2,I}(5247)$, but $\Delta\lambda_V(5250) = 1.23\Delta\lambda_V(5247)$ present a very strong and direct evidence to the kG fields in the solar faculae. As it follows from calculations, more exact value for this case corresponds to $B_f = 1.28 \pm 0.03 \text{ kG}$. This value agrees with magnetic field strengths $1.26\text{--}1.72 \text{ kG}$ obtained by Solanki and Stenflo (1984) from regression coefficients for about 400 unblended FeI lines.

Similar evidences to the strong non-spot field were obtained also with Echelle spectrograph of the horizontal solar telescope of the Kyiv Shevchenko National University (Kurochka *et al.* 1980). In particular, for the places of photosphere in active region AR 19864 of 12 October 1977 the following Stokes V peak ratio was found: $\Delta\lambda_V(5250)/\Delta\lambda_V(5247) = 1.20$. If we add to these data, the necessary data about $\Delta\lambda_{1/2,V}$ and $\Delta\lambda_{1/2,I}$, then we obtain $B_f = 1.41 \pm 0.10 \text{ kG}$.

Similar values of the non-spot fields were obtained with the same instrument using Stokes V separation method for other dates as well. For instance, for the 2B flare of 16 June 1989, the fluxtube magnetic field strength was found to be about 1.1 kG at the start of the flare and during the flash phase, 1.55 kG during the peak, and 1.38 kG 16 min after the peak (Lozitska & Lozitsky 1994). For the 1N solar flare of 26 June 1981, the value of $B_f \approx 1.8 \text{ kG}$ was obtained for the flare peak using the observations in FeI 6301.5 and 6302.5 lines (Lozitsky *et al.* 1999).

During 2000–2004 new observations of the non-spot fields were carried out with named Echelle spectrograph in Kyiv (Kurochka *et al.* 1980), and these observations allow to make the following general conclusion. In each case when the Zeeman splitting is reliably measured on the spectrogram ($B_{\parallel} > 100 \text{ G}$) we found the effects in Stokes V separation of the FeI 5250 and 5247 lines which is similar to the one discovered earlier by Stenflo *et al.* (1987). So, these new data independently confirm earlier observations and strongly indicate to the reality of the spatially unresolved magnetic fields of kG range.

4. Method by Zirin and Cameron

This method was proposed for observations with the Big Bear Spectrovideomagnetograph (Zirin & Cameron 2002). In this instrument, signal of the circular polarization V is integrated on spectral intervals $\pm(0 \div 150) \text{ mÅ}$ relatively the line center, which

for a line like FeI 5250.2 correspond to whole spectral line – from line core to its far wings. Normalized magnetographic signal in line λ_1 , M_{λ_1} , could be presented via formula

$$M_{\lambda_1} = \int_{\lambda_l}^{\lambda_v=0} V/g_{\lambda_1} d\lambda - \int_{\lambda_{V=0}}^{\lambda_r} V/g_{\lambda_1} d\lambda, \quad (3)$$

where first integral is calculated in spectral interval from $-150 \text{ m}\text{\AA}$ to the Stokes V zero crossing point, and second one – from named point to $+150 \text{ m}\text{\AA}$.

After some mathematical transformations, named authors obtained

$$M = 2\alpha_m z \lambda_0^2 B D_m, \quad (4)$$

where α_m is possible filling factor for magnetic component, z – some constant, and D_m – the line depth in the magnetic component for the Stokes I .

Such presented value M was compared for FeI 5247 and 5250 lines, and obtained results are given in Fig. 3 in paper by Zirin & Cameron (2002). From this figure it follows that values M_{5247} and M_{5250} correlate very closely in strength range $-50 \text{ G} < M < 100 \text{ G}$, and $M_{5247}/M_{5250} \approx 1$. So, there was no evidence of the differences of 10–30%, which were discovered earlier by Stenflo (1973) and other authors.

Zirin & Cameron (2002) have considered once more diagnostic dependence (see Fig. 4 in named article): difference $V_{5250}/3 - V_{5247}/2$ versus distance from line center $\Delta\lambda$. Theoretically, for really weak or moderate magnetic field ($< 200\text{--}300 \text{ G}$) this difference should be equal to zero, whereas for strong fields ($> 1 \text{ kG}$) – far from zero. Named authors obtained some non-zero difference from their observations, but it was about three times less than for kilogauss fields.

For a critical analysis of Zirin & Cameron (2002), we should firstly point out that one cannot fully exclude some non-typical case of the difference between FeI 5247 and 5250. For example, in the paper by Cerdena *et al.* (2003) the following line ratio for inter-network magnetic fields was obtained: $B_{\text{eff}}(630.15)/B_{\text{eff}}(630.25) = 1.25 \pm 0.14$. This line ratio indicates the true magnetic field of $B_f = 1.25\text{--}1.5 \text{ kG}$ in the small-scale unresolved fluxtubes. Note this result bases on observations with very high spatial resolution (0.2 arc sec, or 150 km), which were interpreted with a method like line ratio technique. Maybe, for any local places on the Sun the line ratio could be very different from the majority cases. For example, Lozitsky & Lozitsky (2002) discovered a practically opposite ratio for regions of flares: $B_{\text{eff}}(630.15)/B_{\text{eff}}(630.25) = 0.8\text{--}1.1$. Perhaps, similar non-typical (but non-flare) cases are present in the observational statistics in Zirin & Cameron (2002).

On the other hand, we can point to another, systematic source of the alternative conclusion made by these authors. As it was pointed above, in a case of very wide integrating intervals for Stokes V signal, the useful information about small-scale fields should be lost. To preserve this information, an enough high spectral resolution is needed. In fact, as it follows from our calculations, theoretical dependencies for line ratio k (see Fig. 2) have the following peculiarities: $k < 1$ for $\Delta\lambda < 70\text{--}90 \text{ m}\text{\AA}$, but $k > 1$ for $\Delta\lambda > 70\text{--}90 \text{ m}\text{\AA}$. Naturally, we can expect $k \approx 1$ in more wide spectral interval $\Delta\lambda = 0\text{--}150 \text{ m}\text{\AA}$ (as in BBSO instrument). So obtained by Zirin & Cameron's (2002) ratio $M_{5247}/M_{5250} \approx 1$ has a very natural explanation,

which does not exclude the existence of the small-scale magnetic fluxtubes with kG fields.

5. Discussion

It will be useful to remember that the first magnetographic observations by Severny (1967) discovered *non-constant* line-to-line strength ratio for non-spot regions. He obtained the *various* ratio $B_{\parallel}(6103)/B_{\parallel}(5250)$ including the cases when, for example, $B_{\parallel}(6103)/B_{\parallel}(5250) \approx 0$, and $B_{\parallel}(6103)/B_{\parallel}(5250) \approx 1$. Severny (1967) interpreted these results as a manifestation of the very inhomogeneous magnetic field structure in vertical direction.

Later, *constancy* of the line-to-line strength ratio was measured by Harvey & Livingston (1969). They discovered practically constant line ratio $B_{\parallel}(\lambda_i)/B_{\parallel}(\lambda_j)$ from magnetographic measurements in several spectral lines (λ_i and λ_j) including well known lines as FeI 5250 and 5233. For example, the ratios of $B_{\parallel}(5233)/B_{\parallel}(5250) \approx 3$ were found. Exactly, the constancy of these line ratio was measured in a limited field range only, about $B_{\parallel} < 100 - 200$ G. This result was confirmed late by Stenflo (1973) and Gopasyuk *et al.* (1973).

Stenflo (1975) reasonably pointed that constancy of the line ratio presents a very important information about physical characteristics of the spatially unresolved fluxtubes. So long as $B_{\parallel}(\lambda_i)/B_{\parallel}(\lambda_j) \approx \text{const}$, all fluxtubes should have *universal and homogeneous properties*. In particular, all fluxtubes should have the same magnetic field strengths, the same field profiles, the same thermodynamical conditions and Doppler shifts.

Without doubt, this presents an important mystery of the solar magnetism. In fact, during the acquisition of the magnetogram (about several minutes to several tens of minutes) we observe many fluxtubes, which have a relatively short life-time (about several minutes, as it follows from filigree observations) and somewhere occur, somewhere decay. So, we can expect, that their magnetic field strengths and other parameters *are not constant* during the observations. But, in this case, we should observe *no constant* line ratio – for example, from $B_{\parallel}(5233)/B_{\parallel}(5250) = 2-3$ for maximum field concentration to $B_{\parallel}(5233)/B_{\parallel}(5250) \approx 1$ for a time when the fluxtube is born or when it decays. Finally, we must obtain no narrow dependence on “ $B_{\parallel}(5233) - B_{\parallel}(5250)$ diagram”, but a *wide sector of measures*. Wonderfully, we do not observe this effect!

One of possible explanations is the following. Maybe, processes of the birth and decay are very short in time, and escape our attention on ‘line ratio diagram’. On the other hand, new observations by Cerdana *et al.* (2003) indicate *some dispersion* of line ratio, namely $B_{\text{eff}}(630.15)/B_{\text{eff}}(630.25) = 1.25 \pm 0.14$, which could indicate the ‘trace’ of the rapid stages (begin and final) of the fluxtube evolution. Another possibility: perhaps, is that during the fluxtube evolution their main magnetic characteristics are constant, although their diameter changes essentially.

No constant line ratios was obtained earlier by Lozitsky & Tsap (1989). They published the observations for the quiet regions with *two line ratios*, namely $B_{\parallel}(5184)/B_{\parallel}(5253) = 1.50$ and 1.15 . As it follows from model calculations, first case corresponds to fluxtube field, whereas second one – almost homogeneous field like ‘canopy’ (Giovanelly 1980) at Mg I 5184 formation level.

So, the problem of line strength ratio variability is very important, practically unexplored and needs an additional study in the future.

6. Conclusion

Although the kilogauss concept occurred firstly from the line-to-line strength ratio problem, there is a separate problem of the line ratio differences. The reality of the kG fields in solar small-scale fluxtubes follows not only from line ratio method, but also from other techniques: spectral-polarized observations (see e.g., Kouchmy & Stellmacher 1978; Lozitsky 1980, 1986; Lozitska & Lozitsky 1994; Lozitsky *et al.* 2000, etc.), Fourier transformation technique (Stenflo *et al.* 1984; Solanki & Stenflo 1984; Keller *et al.* 1990, etc.), magnetic field measurements using the infrared lines (Harvey & Hall 1975; Stenflo 1989; Muglach & Solanki 1992, etc.), and last observations with best parameters of the direct spatial resolution (Cerdena *et al.* 2003).

From all these observations, it follows that solar fluxtubes with kG fields *occur on the Sun anywhere* – from very quiet regions to active regions and flares, and likely, from center of solar disk to solar pole. As to last, let us remember that about 30 years ago Gopasyuk *et al.* (1973) measured $B_{\parallel}(6103)/B_{\parallel}(5250) = 1.45$ at disk center, and $B_{\parallel}(6103)/B_{\parallel}(5250) \approx 1.3$ at north pole. In the case of homogeneous field, we should obtain $B_{\parallel}(6103)/B_{\parallel}(5250) \approx 1$ in both cases. So, these very old, but very important observations indicate indirectly that fluxtube structure of the solar magnetic field presents on the Sun even outside the zone $|\Delta\phi| = 5\text{--}40^\circ$, where the sunspots occur. It will be very interesting to compare the magnetic properties of fluxtubes in the named zone and outside the zone, but this task needs additional and careful studies in the future.

References

- Babcock, H. W. 1953, *Astrophys. J.*, **118**, 387.
 Cerdena, D., Almeida, J. S., Kneer, F. 2003, *Astron. Astrophys.*, **407**, 741.
 Carroll, T. A., Staude, J. 2001, *Astron. Astrophys.*, **378**, 316.
 Frazier, E. N., Stenflo, J. O. 1978, *Astron. Astrophys.*, **70**, 789.
 Giovanelli, R. G. 1980, *Solar Phys.*, **68**, No. 1, 49.
 Gopasyuk, S. I., Kotov, V. A., Severny, A. B., Tsap, T. T. 1973, *Solar Phys.*, **31**, 307.
 Harvey, J., Hall, D. 1975, *Bull. Amer. Astron. Soc.*, **7**, 459.
 Harvey, J., Livingston, W. 1969, *Solar Phys.*, **10**, 283.
 Harvey, J., Livingston, W., Slaughter, C. 1972, in Proc. Conf. on *Line Formation in Magnetic Field*, NCAR, Boulder, Colorado, p. 227.
 Howard, R., Stenflo, J. O. 1972, *Solar Phys.*, **22**, 402.
 Keller, C. U., Solanki, S. K., Steiner, O., Stenflo, J. O. 1990, *Astron. Astrophys.*, **233**, 583.
 Kotov, V. A., Severny, A. B., Tsap, T. T. 1982, *Izv. Krimsk. Astrofiz. Obs.*, **65**, 3.
 Kouchmy, S., Stellmacher, G. 1978, *Astron. Astrophys.*, **67**, 93.
 Kurochka, E. V., Kurochka, L. N., Lozitsky, V. G., Lozitska, N. I., Ostapenko, V. A., Polupan, P. N., Romanchuk, P. R., Rossada, V. M. 1980, *Vestn. Kiev. Univ., Ser. Astronomii*, **22**, 48.
 Lozitska, N. I., Chesnok, Yu. A., Lozitsky, V. G. 1982, *Vestn. Kiev. Univ., Ser. Astronomii*, **24**, 45.
 Lozitska, N., Lozitsky, V. 1994, *Solar Phys.*, **151**, 319.
 Lozitsky, V. G. 1979, *Astrometria i Astrofizika*, **38**, 13.
 Lozitsky, V. G. 1980, *Phys. Solariterr., Potsdam*, **14**, 88.
 Lozitsky, V. G. 1986, *Kinematic Phys. Celest. Bodies*, **2**, 28.
 Lozitsky, V. G., Dolgoplov, V. I. 1983, *Solnechnyje Dannya*, **5**, 71.
 Lozitsky, V. G., Lozitska, N. I., Lozitsky, V. V., Baranyi, T., Ludmany, A., Mezo, Gy. 1999, Proc. 9th Europ. Meeting on *Solar Physics*. Florence, Italy: ESA Sp-448, 853.
 Lozitsky, V. G., Baranovsky, E. A., Lozitska, N. I., Leiko, U. M. 2000, *Solar Phys.*, **191**, 171.
 Lozitsky, V. G., Lozitsky, V. V. 2002, *Izv. Krimsk. Astrofiz. Obs.*, **98**, 202.
 Lozitsky, V. G., Tsap, T. T. 1989, *Kinematika i Fizika Neb. Tel*, **5**, 50.

- Muglach, K., Solanki, S. K. 1992, *Astron. Astrophys.*, **263**, 301.
- Rachkovsky, D. N. 1963, *Izv. Krivsk. Astrofiz. Obs.*, **29**, 97.
- Rachkovsky, D. N., Tsap, T. T. 1985, *Izv. Krivsk. Astrofiz. Obs.*, **71**, 79.
- Severny, A. B. 1964, *Space Sci. Rev.*, **3**, 451.
- Severny, A. B. 1965, *Astron. Zhurn.*, **42**, 217.
- Severny, A. B. 1966, *Astron. Zhurn.*, **43**, 465.
- Severny, A. B. 1967, *Izv. Krivsk. Astrofiz. Obs.*, **36**, 22.
- Semel, M. 1981, *Astron. Astrophys.*, **97**, 75.
- Solanki, S. K., Stenflo, J. O. 1984, *Astron. Astrophys.*, **140**, 185.
- Staude, J. 1970a, *Solar Phys.*, **12**, 84.
- Staude, J. 1970b, *Solar Phys.*, **15**, 102.
- Staude, J., Hofmann, A. 1988, In: *Proc. Int. Workshop on Reconnection in Space Plasma*, Potsdam; ESA SP-285 Vol. **II**, 123.
- Stenflo, J. O. 1973, *Solar Phys.*, **32**, 41.
- Stenflo, J. O. 1975, *Solar Phys.*, **42**, 79.
- Stenflo, J. O. 1989, *Astron. Astrophys. Rev.*, **1**, 3.
- Stenflo, J. O., Harvey, J. W., Brault, J. W., Solanki, S. 1984, *Astron. Astrophys.*, **131**, 333.
- Stenflo, J. O., Solanki, S., Harvey, J. W. 1987, *Astron. Astrophys.*, **171**, 305.
- Steshenko, N. V. 1967, *Izv. Krivsk. Astrofiz. Obs.*, **37**, 21.
- Wiehr, E. 1978, *Astron. Astrophys.*, **69**, 279.
- Zirin, H., Cameron, R. 2002, *IL Nuovo Cimento*, 25C, No. 5–6, 557.

Erratum

An Apparent Phenomenal Descriptive Method for Judging the Synchronization of Rotation of Binary Stars

Li Lin-sen

(*J. Astrophys. Astr.*, (2004) Vol. 25, Nos. 3 & 4, pp. 203–211)

Table 1. The calculated results for synchronous parameters of twenty components in ten eclipsing binary systems by using the apparent descriptive method of formula (4).

Name		Sp type	P (d)	i (deg)	$R_{1,2}(R_{\odot})$	$(V_{1,2} \sin i)_M$ (km/s)	$Q_{e,e'}$	$P_R \gtrless P$	Synch
GT Cep	1	B2V	4.9087	78	4.64	42	1.11	$P_R > P$	B
	2	A0			6.04	58	1.05	$P_R > P$	A
My Cyg	1	Am	4.0051	89	2.32	26	1.13	$P_R > P$	B
	2	Am			2.06	47	0.55	$P_R < P$	D
V451 Oph	1	B9V	2.1966	87	2.48	48	1.19	$P_R > P$	B
	2	A0			1.98	45	1.01	$P_R > P$	A
U Oph	1	B4.5	1.6773	86.6	3.27	110	0.89	$P_R < P$	B
	2	B5.5			3.02	100	0.91	$P_R < P$	B
V505 Sgr	1	A2v	1.8287	75	2.30	100	0.95	$P_R < P$	A
	2	F81V			2.35	50	1.94	$P_R > P$	E
CV Vel	1	B2V	6.8925	88.2	4.27	50	0.63	$P_R < P$	D
	2	B2V			4.16	85	0.36	$P_R < P$	D
CD Tau	1	F7V	3.4351	88	1.49	34	0.64	$P_R < P$	D
	2	F7V			1.50	34	0.65	$P_R < P$	D
V1143 Cyg	1	F5V	7.6407	87	1.34	9	0.98	$P_R < P$	A
	2	F5V			1.04	20	0.34	$P_R < P$	D
ZZ Boo	1	F2V	4.9917	88	1.76	12	1.48	$P_R > P$	E
	2	F2V			1.70	25	0.69	$P_R < P$	D
AS Cam	1	B8V	3.4309	89	2.06	40	0.7	$P_R < P$	C
	2	A0V			2.47	30	1.21	$P_R > P$	C

Table 2. The calculated results for synchronous parameters of thirty-five components in twenty single-lined and double-lined spectroscopic binary systems by using four apparent descriptive methods.

The first method calculated by using $a_{1,2} \sin i$										
Name	Sp type	P (d)	$R_{1,2}$ ($R_{\odot}$)	$m_{1,2}$ ($m_{\odot}$)	$q_{1,2}$	$(V_{1,2} \sin i)_M$ (km/s)	$(a_{1,2} \sin i) \times 10^6$ (km)	$Q_{1,1'}$	$P_R \leq P$	Synch
AG Per	1 B5pV	2.0287	3.02	5.08	0.89	105	4.53	0.69	$P_R < P$	D
	2 F6		2.78	4.52	1.1	120	4.97	0.55	$P_R < P$	D
V448 Cyg	1 O9.5	6.5179	10.00	22.25	0.73	95	$1.92 \times 10^{+1}$	1.08	$P_R > P$	B
	2 BII		17.34	16.24	1.37	92	$1.50 \times 10^{+1}$	1.10	$P_R > P$	B
GK Cep	1 A2V	0.9362	2.67	2.67	0.93	106	2.21	1.29	$P_R > P$	C
	2 A4		2.45	2.48	1.08	100	2.41	1.27	$P_R > P$	C
EI Cep	1 FI	8.4397	2.99	1.80	0.95	15	9.42	1.25	$P_R > P$	C
	2 FI		2.31	1.71	1.05	10	8.92	1.31	$P_R > P$	D
AH Cep	1 O8	1.7748	6.44	16.52	0.87	175	6.08	1.08	$P_R > P$	B
	2 O9		5.98	14.37	1.15	160	6.91	1.02	$P_R > P$	A
The second method calculated by using $m_{1,2} \sin^3 i$										
Name	Sp type	P (d)	$R_{1,2}$ ($R_{\odot}$)	$m_{1,2}$ ($m_{\odot}$)	$q_{1,2}$	$(V_{1,2} \sin i)_M$ (km/s)	$m_{1,2} \sin^3 i$ ($m_{\odot}$)	$Q_{2,2'}$	$P_R \leq P$	Synch
EK Cep	1 A1V	4.4278	1.59	2.11	0.55	45	2.0	0.63	$P_R < P$	D
	2 G5		1.19	1.16	1.82	14	1.1	1.02	$P_R > P$	B
NY Cep	1 B01V	15.2767	10.56	16.51	0.67	80	$2.7 \times 10^{+1}$	0.52	$P_R < P$	D
	2 B01V		8.66	11.06	1.49	71	$1.3 \times 10^{+1}$	0.43	$P_R < P$	D
RX Her	1 B9.5 V	1.7786	2.37	2.86	0.86	73	2.7	0.93	$P_R < P$	B
	2 A0.5 V		2.05	2.50	1.07	60	2.3	0.94	$P_R < P$	B
Z Vul	1 B4V	2.4454	4.54	5.42	0.43	98	5.4	0.95	$P_R < P$	A
	2 A3III		4.57	2.33	2.32	115	2.2	0.79	$P_R < P$	C
λ Tau	1 B3V	3.9529	5.64	6.04	0.24	85	6.4	0.87	$P_R < P$	B
	2 A4IV		4.33	1.45	4.16	60	1.7	0.97	$P_R < P$	A

Table 2. (Continued)

Name		Sp type	$P(d)$	The third method calculated by using $K_{1,2}$ ($V_{1,2} \sin i)_M$				$Q_{3,3'}$	$P_R \leq P$	Synch
				$R_{1,2}$ ($R_\odot$)	$m_{1,2}$ ($m_\odot$)	$q_{1,2}$	e			
EG Ser	1	A0	9.9473	2.17	3.05	0.78	0	0.22	$P_R < P$	D
	2	A2		1.93	2.37	1.29	0	0.12	$P_R < P$	D
V624 Her	1	A _{3m}	3.8950	2.93	2.23	0.86	0	1.02	$P_R > P$	A
	2	A _{4m}		2.32	1.90	1.17	0	0.84	$P_R < P$	B
Sig Aql	1	B ₃ V	1.9503	4.12	6.48	0.86	0	0.90	$P_R < P$	B
	2	B ₃ V		3.44	5.58	1.16	0	0.71	$P_R < P$	C
WW Air	1	A7V	2.5250	1.93	1.81	0.96	0	1.09	$P_R > P$	B
	2	A7V		1.92	1.74	1.04	0	1.15	$P_R > P$	B
RZ Cha	1	F5IV	2.8321	2.94	1.90	0.84	0	1.34	$P_R > P$	D
	2	F5IV		3.19	1.59	1.19	0	1.27	$P_R > P$	C
Name		Sp type	$P(d)$	The fourth method calculated by using $f(m)$ ($V_1 \sin i)_M$				Q_4	$P_R \leq P$	Synch
				R_1 ($R_\odot$)	m_1 ($m_\odot$)	q_1	$f(m)$ ($m_\odot$)			
UW Vir ₁		A ₄	1.8107	1.52	1.77	0.24	76	1.9×10^{-2}	$P_R < P$	D
TW Dra ₁		A ₅ V	2.8067	2.06	2.24	0.43	123	5.9×10^{-2}	$P_R < P$	D
QS Aql ₁		B5V	2.4968	2.65	6.82	0.17	75	2.7×10^{-2}	$P_R < P$	C
DV Aqr ₁		LateA	1.5755	2.16	2.20	0.59	92	1.4×10^{-1}	$P_R < P$	C
UX Her ₁		A3	1.5489	1.80	1.92	0.30	61	2.1×10^{-2}	$P_R < P$	B

References

Li, Lin-sen, *Acta Astrophys. Sinica*, 1998, **18**, 77; 1997, **17**, 407.