

Stellar and Extragalactic Radiation at the Earth's Surface

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Abstract. Reviving a calculation made by Eddington in the 1920s, and using the most recent and comprehensive databases available on stars and galaxies, including more than 2,500,000 stars and around 20,000 galaxies we have computed their total radiation received at the Earth just outside its atmosphere. This radiation density, if thermalized, would be equivalent to a temperature of 4.212 K. The comparability of this temperature to that of the cosmic microwave background (2.723 K) may either be a pure coincidence or may hold a key to some as yet unknown, aspect of the universe.

Key words. Stellar background—stars—CMBR.

1. Introduction

In 1926, Eddington made an order of magnitude calculation in which he estimated the total radiation background from stars in the Galaxy by assuming a population of only 2000 stars of apparent bolometric magnitude $m = 1.0$. He arrived at an energy density of starlight of around 7.67×10^{-13} erg cm⁻³. If equated to a black body equilibrium distribution, this worked out to a temperature of ~ 3 K. Eddington identified it with the 'temperature of interstellar space', agreeing with an earlier calculation of Ch. Fabry that this was certainly an underestimation.

Today we have much better and far more comprehensive databases than those available to Eddington. Using them it is possible to redo the calculation which typically runs along the following lines.

2. The basic equations

The total radiation received from all directions by a square centimetre of surface area per second at the Earth will be denoted by

$$j = j_s + j_g \quad (1)$$

of which j_s is the radiation of N stars in the Galaxy (which also includes those in the Milky Way band) and j_g is the contribution of P external galaxies. The Sun is excluded from the first term and our Galaxy from the second.

To fix units we shall assume that a light source of *zero bolometric magnitude* ($m_{\text{bol}} = 0$) gives an energy flux j_0 in the vicinity of the Earth, where

$$j_0 = 2.48 \times 10^{-5} \text{ erg cm}^{-2} \text{ s}^{-1} \quad (2)$$

as given by Allen (1999). Thus we can write the corresponding energy flux from a light source of bolometric magnitude m as

$$j(m) = j_0 \times 10^{-0.4m} = j_0 \times \text{dex}(-0.4m). \quad (3)$$

Assuming that we have a stellar distribution in which there are n_i stars of magnitude m_i , with $i = 1, 2, 3, \dots$, we get

$$N = \sum_i n_i, \quad (4)$$

$$j_s = j_0 \sum_i n_i \times \text{dex}(-0.4m_i). \quad (5)$$

Likewise for galaxies we shall write

$$P = \sum_i p_i, \quad (6)$$

$$j_g = j_0 \sum_i p_i \times \text{dex}(-0.4m_i). \quad (7)$$

We will use these formulae in estimating the stellar and extragalactic contributions.

3. Stellar radiation

We have used the data given by the CDS, Strasbourg for $N = 2,552,323$ stars. The data are available in tables grouping stars in magnitude intervals of $\Delta m = 0.25$, either in BT or VT magnitudes. To be more rigorous and consistent, one should use only *one* magnitude (either B or V) and correct it properly for the bolometric correction (BC) derived from the colour indices.

However, there are some caveats. Firstly, we are ignoring the high frequency (UV, X, γ) radiation from active stars; this procedure therefore underestimates their radiative contributions. These are, however, negligible for our purpose here. We also note that most of the radiation lost by reddened stars is converted to heat or to excite the interstellar matter and is thus lost to us.

Secondly, we are using a catalogue that may not be complete for faint stars. Therefore while applying equation (5) we may have to correct the apparent luminosity function by a suitable extension to large magnitudes. This effect is, however, small and by not including it we may underestimate j_s by a very small amount.

Thirdly, the data list stars binned in intervals of 0.25 magnitude, with data on all stars brighter than $m = 0$ having been put together. This is corrected for in our calculation by taking data on individual stars with magnitudes up to $m = 2.5$. We do not expect different binning to produce a significantly different answer since the number-magnitude distribution is a fairly smooth function.

While calculating j_s from the tables, one needs to use a characteristic average value of m_i for the i th bin. The shape of the number-magnitude $n(m)$ curve helps in settling this issue. The curve for B-magnitudes has a linear rising trend for $4 < m < 11$, and

a linear descending trend for $m > 14$. One may therefore write for each of the above ranges,

$$\Delta \log n(m) = k \Delta m, \quad k = \text{constant}, \quad (8)$$

where Δ relates to each typical interval ($m, m + 0.25$). The average for this interval is then given by

$$\langle m \rangle = k^{-1} \log \left[\frac{\{\text{dex } km + \text{dex } k(m + 0.25)\}}{2} \right]. \quad (9)$$

To deal with the rather imprecise CDS statistics for bright stars, we have taken them individually to form a sample of 72 stars with magnitudes $m < 2.5$. We find that the total stellar contribution to the background is made up of set A: faint or moderately faint stars with $m > 2.5$ and set B: bright stars with $m < 2.5$. These add up to:

$$j_s = j_{sA} + j_{sB} = 1.782 \times 10^{-2} \text{ erg cm}^{-2} \text{ s}^{-1}. \quad (10)$$

4. Radiation from galaxies

Before considering external galaxies, it is necessary to note that among earlier authors, Pannekoek *et al.* (1949) had estimated the Milky Way contribution to the night sky background using out of focus photographic photometry. Although this value is not quoted by standard sources like Allen (1999), we give it here as a curiosity after correcting for BC:

$$j_{MW} = 1.529 \times 10^{-2} \text{ erg cm}^{-2} \text{ s}^{-1}. \quad (11)$$

Remarkably, this was close to our estimate (equation 10) obtained by stellar counts. We shall therefore assume that the Milky Way contribution is implicitly incorporated in equation (10) and proceed with this value in what follows.

For external galaxies we have used the catalogue of de Vaucouleurs *et al.* (1991). Dividing the 24 hours of right ascension into 24 sets we have used subsets of 10-minute interval each as representative samples. Then we have multiplied the figures for each magnitude interval by 6 and obtained an estimate of $p(m)$, the number of galaxies in a typical magnitude interval ($m, m + 1$). A method similar to that used for stars but with $\Delta m = 1$, instead of 0.25, is used to estimate the average magnitude $\langle m \rangle$ for each interval. We have in all 19,822 galaxies in the magnitude range $m > 9$. We denote this set as set C. Additionally, there are bright galaxies with $m < 9$. We refer to the set of such galaxies with magnitudes in the range $9 < m < 20$ as set D, and we have taken the photographic magnitudes for these 15 galaxies as are given by Landolt and Börnstein (1965). Using these magnitudes requires a small bolometric correction, which we have ignored.

The total contribution from external galaxies is then

$$j_g = j_{gC} + j_{gD} = 3.467 \times 10^{-5} \text{ erg cm}^{-2} \text{ s}^{-1}. \quad (12)$$

Thus the total contribution from external galaxies is quite negligible compared to that from stars in the Milky Way.

5. Discussion of the errors

The first source of uncertainty which one can think of, is the existence of variable stars. As one has certainly counted them at their brightest phase, their contribution indicates an overestimation of the energy. However, it must be a very small error. According to our estimation, only about 10^5 variable stars exist in the galactic bulge of $> 10^{10}$ stars. Dropping them completely from the statistics would lead to an error on the energy of less than 10^{-5} . Another way to estimate more generally the number of variable stars is to evaluate the time which they spend as variable in their evolution, from 'birth' to extinction (as white dwarf or neutron star), during all their expected life time. Again, even for a star like the Sun, the first phase of its life (contraction from interstellar space) is of the order of 10^5 years against its stable life of about $> 10^{10}$ years. Its final stages of evolution towards the white dwarf stage is also very short (although one knows little about the time element). Evolution of the Sun might not be indicative of the more variable of the stars; but the variable star phase does not affect all stars; moreover it occupies a very little space in the colour-luminosity diagram of all observable stars, where the largest majority of observable stars are in their phase of stability, on the main sequence. The conclusion of our inquiry is therefore that variability of stars does not affect the energy received at the Earth by more than a factor 10^{-5} .

The existence of supernovae is also extremely small and hardly enough to alter our results. There is at the most 1 supernova per normal galaxy (such as ours) in 50 years; and it lasts schematically only a few months. Assuming a typical luminosity at its maximum of 10^{40} erg s $^{-1}$, assuming for it an average distance of 1 kpc, its energy flux at the Earth is 10^{-4} erg s $^{-1}$ cm $^{-2}$, i.e., 100 times smaller than our estimates of the stellar energy flux at the Earth, and as this happens only during 3 months for every 50 years, this (statistically) diminishes its importance by another factor 10^2 . Of course, supernovae may appear in other galaxies. The case of SN 1987A in the Small Magellanic Cloud is well-known: At a distance of 55 kpc, the flux at maximum phase was only 4×10^{-8} ergs s $^{-1}$ cm $^{-2}$, i.e., 10^{-6} that of our estimate of the stellar flux at the Earth. Our conclusion here is that neglecting the SN in other galaxies affects the results at the 6th significant figure, and neglecting the SN of our own Galaxy affects them at their 4th significant figure. It is difficult to do a better estimation of the error, due to the random character of supernovae explosions; but it seems clear that neglecting them is perfectly allowed, and that our results are essentially not modified by this source of error.

The high energy radiation from the gamma rays, X-rays, and UV part of the spectra, is considered implicitly for 'normal' stars of a quasi-thermal behaviour. But there are strong UV, X-ray and gamma-ray emitters. Are they affecting our results? X-ray sources have been observed by several satellites. Estimating their number as $N = 105$, with a flux at the Earth of 10^{-11} erg s $^{-1}$ cm $^{-2}$ gives us a flux of the order of 10^{-6} erg s $^{-1}$ cm $^{-2}$. About 100 gamma-ray sources have a still much smaller flux (10^{-4} MeV s $^{-1}$ cm $^{-2}$ per source) which amounts to 1.6×10^{-8} erg s $^{-1}$ cm $^{-2}$. There is no estimate of UV sources that would not be either normal stars, or X-ray or gamma-ray sources. Altogether, the figures above indicate again that only the fourth significant figure of our estimate of the stellar flux at the Earth could thus be affected.

Nevertheless, the estimation we have made concerns only the present state of stellar and extragalactic radiation received **now** at the Earth. Clearly, this received radiation certainly varies during the life span of our Galaxy, even without considering the

changes due to cosmological evolution of the environment of our Galaxy. This point becomes important as soon as we try to understand the background radiation of the Earth's sky.

6. Equilibrium temperature

Assuming that this radiation background were somehow thermalized, what black body temperature T would it produce? The answer is simple and is obtained by using the Stefan–Boltzmann law

$$\sigma T^4 = j_s + j_g. \quad (13)$$

For the value given by the equation (10) we find that $T = 4.212$ K. If we consider, however, a typical point in the intergalactic medium, far from our Galaxy, we need to use j_g only, as in equation (12) and get $T = 0.88$ K.

The comparability of the first value, 4.212 K to the temperature of the cosmic microwave background (CMB), namely 2.723 K makes one wonder if this is a pure coincidence. Perhaps this suggests a possible physical connection between starlight energy density and the energy density of CMB. It has already been noticed (for a recent discussion, see Hoyle *et al.* 2000) that if all of cosmic helium were processed in stars, the resulting starlight energy density would come very close to that of the CMB.

A cosmologist would point out that at a general point in the universe the coincidence of temperatures disappears and so any link with CMB is spurious. This is a possible position to take and it will certainly be taken in standard cosmology. Nevertheless, we feel that this modern reworking of Eddington's estimate of radiation background is of interest and may hold a key towards a deeper understanding of radiation backgrounds in the universe.

7. Conclusion and further research

Needless to say, our simple-minded result confirms that of Eddington's. It is important to note that we thus found a value of the equilibrium temperature quite of the same order of magnitude as that of the measured radiation of the microwave background, only slightly hotter. This rather good coincidence may be of a great significance.

However, before attempting any identification between our computations and the available observations, we have

- to compute, in the solar system, what can be the time scale of thermalisation in a very transparent medium,
- to explain why in this case, the thermalisation is not quite complete,
- to understand why the bipolar component could thus be explained, and
- to propose an interpretation of the fluctuations now attributed to fluctuations of the microwave background. We shall try to reply to these questions in the future.

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in all colours and all locations on the sky. We have not used the wealth of these data to their full extent, but we hope to use them to refine the present results.

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The Tully–Fisher Relations of the Eridanus Group of Galaxies

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Abstract. The Tully–Fisher (TF) or the luminosity–linewidth relations of the galaxies in the Eridanus group are constructed using the HI rotation curves and the luminosities in the optical and in the near-infrared bands. The slopes of the TF relations (absolute magnitude vs $\log 2V_{\text{flat}}$) are -8.6 ± 1.1 , -10.0 ± 1.5 , -10.7 ± 2.1 , and -9.7 ± 1.3 in the R, J, H, and K bands respectively for galaxies having flat HI rotation curves. These values of the slopes are consistent with those obtained from studies of other groups and clusters. The scatter in the TF relations is in the range 0.5–1.1 mag in different bands. This scatter is considerably larger compared to those observed in other groups and clusters. It is suggested that the larger scatter in the TF relations for the Eridanus group is related to the loose structure of the group. If the TF relations are constructed using the baryonic mass (stellar + HI + Helium mass) instead of the stellar luminosity, nearly identical slopes are obtained in the R and in the near-infrared bands. The baryonic TF (baryonic mass vs $\log 2V_{\text{flat}}$) slope is in the range 3.5–4.1.

Key words. Galaxy; rotation, luminosity–linewidth relation — galaxies: groups — individual: Eridanus — radio lines: HI 21 cm-line.

1. Introduction

A correlation between the total mass derived using the HI linewidths and optical luminosity was noticed in the late 70 s and early 80 s by a number of investigators (e.g., Roberts 1969; Rogstad & Shostak 1972; Balkowski *et al.* 1974; Shostak 1975). Tully & Fisher (1977) suggested a luminosity–linewidth relation of the form $L \propto V_{\text{rot}}^\alpha$ to measure distances to spiral galaxies. Subsequently, the luminosity–linewidth relation was also known as the Tully–Fisher (TF) relation. Tully & Fisher (1977) demonstrated that using some nearby spiral galaxies for which the distances are known accurately, the luminosity–linewidth relation can be calibrated to infer distances to clusters of galaxies as well as to obtain a value for the Hubble constant. The value of index α has been found to be ~ 3 in the optical bands and ~ 4 in the near-infrared bands.

The Tully–Fisher relation can possibly be best attributed to a tight correlation between dark matter and the total baryonic matter embedded in it. This is because the constant rotation velocity in the outer regions of galaxies is believed to be due to dark matter, and the stellar luminosity directly correlates with the stellar mass, which is the dominant component of the baryonic mass in a galaxy. The cold dark matter models of galaxy formation predict a relation ($M_{\text{baryon}} \sim V_{\text{rot}}^\alpha$) between the baryonic mass M_{baryon} and the rotation velocity V_{rot} , where $\alpha \sim 3$ for non-baryonic dark matter (e.g., van den Bosch 2000; Navarro & Steinmetz 2000) and $\alpha \sim 3.5$ for the collisional baryonic dark matter (Walker 1999). The stellar component of the baryonic mass can be estimated from the stellar luminosity, provided, the stellar mass to light ratio (M^*/L) is known. Bell & de Jong (2001) estimated M^*/L as a function of colour using the spectro-photometric stellar population synthesis models of galaxy evolution. Observationally, the value of α is found in the range 3.5–4.0 (e.g., McGaugh *et al.* 2000; Bell & de Jong 2001). It is advantageous to study the TF relation using the near-infrared luminosities due to the following two reasons:

- the extinction due to dust is minimal in these bands and,
- the total stellar mass in galaxies is dominated by the old stellar population, which emits mostly in the near-infrared.

The I-band TF relation in several groups and clusters has been studied by Giovanelli *et al.* (1997) using HI linewidths and optical rotation curves. They noticed that the TF relations in the Eridanus and Centaurus groups, and in the Fornax cluster have larger scatter ($\sigma \sim 0.32$ – 0.45 mag) compared to that in other groups and clusters ($\sigma < 0.3$). One of the implications of the larger scatter could be significant distance variations for galaxies in those groups or clusters. Giovanelli *et al.* (1997) discussed the selection effects and biases inherent in a sample of galaxies and their consequences on the slope and on the scatter in the TF relations. Furthermore, the observed scatter will also depend on the sample size. In addition to the above effects, various extinction corrections applied to obtain the intrinsic luminosity of a galaxy are not well understood. The linewidth or the rotation velocity also needs to be corrected for the random and turbulent motions in the galactic-disks. Tully & Fouque (1985) have described the procedures to make various corrections to the two TF observables, rotation velocity and the stellar luminosity. Many of these corrections are empirically determined. There are indications that the TF slope depends upon the types of galaxies (Rubin *et al.* 1985). Persic *et al.* (1996) have shown a strong dependence of the shape of the rotation curve on the I-band luminosity. They noticed that galaxies with $M_I < -22.0$ have rotation curves declining in their outer regions. The dependence of the TF relation on the shapes of rotation curves has been recently brought to attention by Verheijen (2001). For galaxies in the Ursa–Major group, they reported a scatter as low as 0.21 mag in the K-band TF relation, consistent with the measurement uncertainties and leaving no scope for any intrinsic scatter in the TF relation. They also reported that the scatter in the near-infrared TF relation is reduced by ~ 0.04 mag if the flat rotation velocity is used instead of the maximum rotation velocity. Therefore, it appears that the flat part of the rotation curve, and hence the dark matter is better correlated with the stellar luminosity.

The TF relations for the Eridanus group have been studied in the I-band (Giovanelli *et al.* 1997) and in the near-infrared bands (Bamford 2002). The I-band TF slope for the Eridanus group (-7.88 ± 0.56) is slightly steeper than the mean slope (-7.6) for

all the clusters studied by Giovanelli *et al.* (1997). Bamford (2002) reported a K-band TF slope of -10.8 and a scatter of 0.58 for the Eridanus group.

In this paper, the optical R-band and the near-infrared TF relations are constructed using the HI rotation curves of galaxies in the Eridanus group using the GMRT observations. The HI observations and the data analyses are described in Omar & Dwarakanath (2005). The near-infrared photometry in the J, H, and K bands are from the *Two Micron All Sky Survey* (2MASS). The optical R-band photometric data are from the 104-cm Sampurnanand telescope at Nainital (Omar 2004).

2. The radio observations

The HI 21 cm-line observations of the Eridanus galaxies were carried out with the Giant Meterwave Radio Telescope (GMRT). The details of the observations and of the data analyses can be found in Omar & Dwarakanath (2005). Here, only a brief description is given. Galaxies were not selected based on their HI content or the optical luminosity. The disk galaxies were selected primarily from the inner 4 Mpc region of the group where the galaxy density is high. This selection criterion was adopted since one of the aims of this survey was to study the galaxy evolution in the high galaxy density regions. After editing and calibrating the data, a channel rms of $\sim 1 \text{ mJy beam}^{-1}$ was obtained. The velocity resolution was typically $\sim 13.4 \text{ km s}^{-1}$ and the spatial resolution was $\sim 20''$ ($\sim 2 \text{ kpc}$).

The velocity fields of galaxies were obtained from one-component Gaussian fittings of the line-profiles at each pixel containing HI signal. It was found that this procedure of constructing velocity field provides better image fidelity compared to that obtained from the moment analysis. The kinematical parameters (position angle, inclination and rotation curve) of the HI disks were obtained using the tilted-ring model. Detailed analyses, maps and results are presented in Omar & Dwarakanath (2005). The galaxies often showed kinematical lopsidedness. However, the rotation curves used in the present analyses are averages of those for the approaching and the receding sides. The rotation curves of the galaxies in the current study have no appreciable difference between the maximum and the flat rotation velocities. Therefore, the TF studies are carried out using the flat part of the rotation curves, which do not require any significant correction for the effects of beam smearing. The flat parts of the rotation curves were inferred by visual inspection.

3. The Optical data

3.1 R-band observations and data reduction

The optical observations were carried out in the R-band (*Cousins*) using the 104-cm Sampurnanand reflector at the Aryabhata Research Institute of observational-scienceS (ARIES, *formerly State Observatory*; longitude $79^\circ 27'$ E, latitude $29^\circ 22'$ N, altitude 1955 m), Nainital. This telescope uses the Ritchey–Chrétien system with a f/13 Cassegrain focus having a plate scale of $15''.5 \text{ mm}^{-1}$. The observations were carried out during October 11–17, 2002. These observations were carried out during a month when the atmospheric conditions at the site are good. The data were recorded on 2048×2048 pixels CCD camera cooled at the temperature of liquid nitrogen. The

Table 1. R-band observations.

Name	Frames	Time	Airmass	FWHM
ESO 482-G 035	2	10	1.71	2.8
ESO 548-G 065	2	10	1.57	2.6
ESO 549-G 002	2	10	1.95	2.8
IC 1952	3	10	1.66	2.3
IC 1953	4	20	1.87	2.9
IC 1962	3	15	1.87	2.4
MCG-03-10-041	2	10	1.76	3.1
NGC 1325	2	10	1.66	2.6
NGC 1371	3	15	1.66	3.2
NGC 1385	6	30	1.65	2.7
NGC 1414	2	10	1.68	2.9
NGC 1422	2	10	1.68	2.9
SGC 0321.2-1929	2	30	1.54	2.0
UGCA 068	2	10	1.77	3.6
UGCA 077	2	10	1.59	3.1

Notes: **Column 1:** Name of the galaxy. **Column 2:** Number of frames. **Column 3:** Total exposure time (in minutes) combining all the frames. **Column 4:** The mean value of the airmass estimated at the mid-exposure time of the frames. **Column 5:** The FWHM (in arcsec) of stellar images on CCD estimated after co-adding the frames.

CCD covers a field of view $13' \times 13'$ on the sky with each square element having a resolution of $0''.38 \times 0''.38$. The data were acquired after averaging the counts in every 2×2 pixels making each square element equal to $0''.76 \times 0''.76$. The read-out noise and the gain of the CCD camera are $10 e^-$ and $5.3 e^-$ per analog-to-digital count respectively. Two to five frames each of 5 min or 3 min duration were observed for each galaxy. A few galaxies having low surface brightness features were observed for a total of 30 min duration. The observational parameters are given in Table 1.

On October 13, 2002 the standard Landolt field SA 92 (Landolt 1992) was observed in the V and R bands for photometric calibration and local atmospheric extinction correction. A total of eight stars from this Landolt field with V magnitudes between 12.5 and 15.6, and, (V–R) colours between 0.31 and 0.72 were observed at different zenith angles. The value of the zero point was 20.79 ± 0.03 and that of the extinction coefficient (k) was 0.18 ± 0.02 on that night. The value of the colour coefficient to make the colour correction was insignificant. Hence, no colour corrections were made. The typical value for the R-band sky brightness was $20.5 \text{ mag arcsec}^{-2}$. The average $FWHM$ was $2''.8 \pm 0''.5$. Based on the vast observing experience at the observatory, if weather is stable during the observing period, no significant variations in the calibration is expected during the nights. The weather during the observing period was found to be stable. Hence, the Landolt field to determine the calibration was observed during one of the five nights. The calibration errors are expected to be less than 0.05 mag, estimated from the fitting errors in the extinction coefficient and the zero-point. These observations lie in the regime where noise in the image is dominated by the sky brightness. We estimate, for a typical galaxy size of $4'$, an additional error of 0.03, 0.15, and 0.7 mag for a 10, 12, and 14 mag galaxy respectively due to Poisson noise.

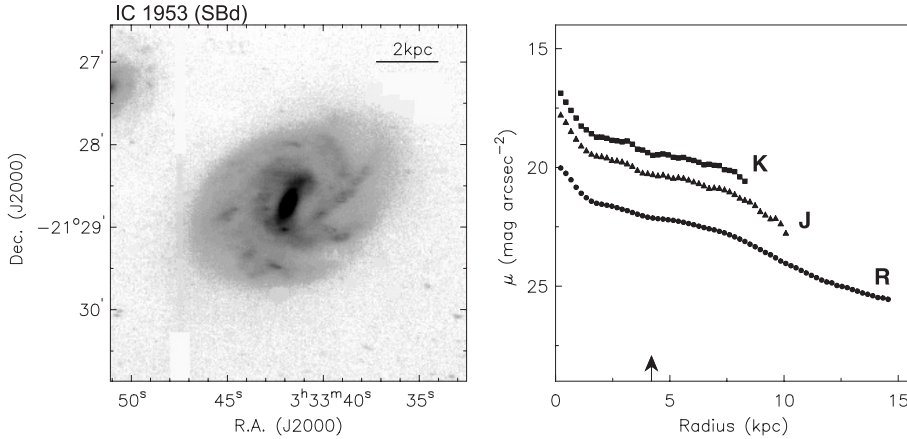


Figure 1. (Left) The R-band image of IC 1953 obtained using the 104-cm telescope at Nainital. (Right) The average surface brightness profiles in R, J, and K bands as a function of galacto-centric distance. The J and K band images were obtained from the 2MASS archival data. The arrow on the x-axis marks the disk scale-length in the R band.

On each night of observation, several bias frames were taken and averaged to correct the CCD frames. Since the exposure time in each frame was short, no significant dark counts are expected. Every alternate night, morning twilight flat fields were observed. For the nights when no flat fields were taken, previous day's flat fields were used. The CCD data reduction was performed using the package CCDRED in the *Image Reduction and Analysis Facility* (IRAF) software package developed by National Optical Astronomy Observatory. The frames were aligned using bright stars in the field. The pixels hit by cosmic rays were removed using the median filtering in each pixel of different frames whenever at least three frames for a field were available. When there were only two frames available, the bad pixels were identified by comparing their values with the average value of the surrounding pixels. After removing the cosmic ray events, all the aligned frames for a galaxy were averaged to get the final image. The images were regridded to a pixel size of $1''$. The images were registered for the equatorial coordinates in the J2000 epoch using the *Digitized Sky Survey* (DSS) plates.

The sky background was taken as the mode of the pixel values in the averaged frame. If the presence of a bright star near the edge or within the observed frame caused significant light gradients across the frame, or, the electron pickup noise caused bright and dark stripes in the frames, a mean value of the sky was estimated using pixels surrounding the galaxy. The Galactic foreground stars over a galaxy were replaced by an average value of the intensities surrounding the affected region. The azimuthally averaged limiting surface brightness is $\sim 26.0 \text{ mag arcsec}^{-2}$ in the R band.

The R-band image of IC 1953 and its surface brightness profile are shown in Fig. 1 as an example. More images can be found in Omar (2004).

3.2 2MASS archival data

The 2MASS is an all sky survey in the near-infrared J ($1.24 \mu\text{m}$), H ($1.66 \mu\text{m}$) and K_s ($2.16 \mu\text{m}$) bands (Jarrett *et al.* 2000). Each 2MASS camera consists of a liquid nitrogen cooled cryostat, housing three 256×256 NICMOS3 CCDs each covering a

field of view $8'.5 \times 8'.5$ on the sky. The final calibrated, co-added (6.8 s integration time), and full resolution images with $1''.0 \times 1''.0$ sampling are obtained from the 2MASS archive.

The overall photometric accuracy of the 2MASS images is believed to be better than ~ 0.1 mag (Jarrett *et al.* 2000). The zero points are typically 20.9, 20.5, and 19.9 for the J, H, and K bands respectively. The sky brightnesses are typically 15.3–16.0, 14.3, 12.9–13.5 mag arcsec $^{-2}$ in the J, H, and K bands respectively. The azimuthally averaged surface brightness sensitivities (3σ) are ~ 21.0 , ~ 20.5 , and ~ 20.0 mag arcsec $^{-2}$ in the J, H, and K bands respectively.

4. Image analyses

The image analyses were carried out using IRAF, KARMA, and GIPSY (*Groningen Image Processing System*). All the images were converted to an identical format with $1''.0 \times 1''.0$ pixel size and regridded to the J2000 epoch. Foreground stars on galaxies, bad pixels, and other defects in the images were removed and replaced by the mean of the pixels surrounding the affected region. The axial parameters of the stellar disks (position angle and ellipticity), surface brightness profiles, colour profiles, disk scale lengths, central disk surface brightnesses, isophotal and total magnitudes, etc. were estimated. Details of the analyses and results are described in Omar (2004). Here, the procedure to obtain the magnitudes of galaxies is described. The photometric data are presented in Table 5 (Appendix 1).

4.1 Isophotal and total magnitudes

Both isophotal and total magnitudes were obtained for the galaxies. The isophotal magnitudes were estimated within the isophote at 26.0, 21.0, 20.5, and 20.0 mag arcsec $^{-2}$ for the R, J, H, and K bands respectively. These are the limiting isophotes for the respective bands. The isophotal magnitude m_l^{iso} corrected for the atmospheric extinction is estimated using the following relation:

$$m_l^{\text{iso}} = Z - 2.5 \log(S_l/t) - k \cdot X, \quad (1)$$

where the subscript l refers to the limiting isophote in a band and the superscript ‘iso’ refers to isophotal, Z is the zero point magnitude corresponding to one second of exposure time, k is the extinction coefficient, X is the airmass, S_l is the total number of counts above the sky background within the limiting isophote, and t is the total exposure time on the galaxy.

The isophotal magnitude provides a lower limit to the actual luminosity of a galaxy. Normally, optical observations are sensitive to detect light up to 4–5 disk scale lengths while near-infrared observations detect light only up to 2–3 disk scale lengths in a reasonable exposure time. It is obvious that different bands are sensitive to different levels in detecting light from galaxies. To overcome this effect, correction for the undetected light has to be applied to obtain the total magnitudes of galaxies. If the disks were infinitely extended and follow the exponential light profiles, it can be shown that the total light above a certain limiting isophotal radius depends only on the number of scale lengths within that limiting radius, and can be estimated using the following relation:

$$\delta m_{\text{ext}} = 2.5 \log(1 - (1 + \delta n) \exp(-\delta n)) \quad (2)$$

(Tully *et al.* 1996); where δm_{ext} is the total light beyond the limiting isophote radius and δn is the number of observed scale lengths within the limiting isophote radius ($\delta n = (\mu_0 - \mu_{\text{lim}})/1.086$). Here, μ_0 and μ_{lim} are the central disk surface brightness and the surface brightness of the limiting isophote respectively. It is suggested by Tully *et al.* (1996) that the disks are in reality not infinitely extended but rather truncated. They empirically determined that only 88% of the δm_{ext} given by equation (2) should be used to make correction to the isophotal magnitude in order to get the total magnitude.

The scale length r_d , and the central disk surface brightness μ_0 were estimated from a fit of the form $I(r) = I(0)e^{-r/r_d}$ to the averaged radial surface brightness profile (see Omar 2004 for details). The average radial surface brightness ($I(r)$) was computed for galaxies in elliptical annuli of width $2''$ with a fixed position angle and ellipticity. The inner regions where the disk profiles are significantly modified due to the presence of a bar or a bulge were excluded in the fit. The range of data to be excluded in the fit was determined by visual inspection.

4.2 Extinction corrections

The total magnitudes were corrected for the Galactic and the internal dust extinction in galaxies. The values of the Galactic extinction provided by Schlegel *et al.* (1998) were used in the present analysis. The high Galactic latitude ($\sim -52^\circ$) of the Eridanus group results in minimal Galactic extinction. The average values of the Galactic extinction toward Eridanus are ~ 0.11 , ~ 0.04 , ~ 0.02 , and ~ 0.01 mag in the R, J, H, and K bands respectively. The estimation of the internal dust extinction in galaxies is quite uncertain. It depends on their orientation, morphological type, size, and luminosity. Several authors have discussed internal dust extinction in galaxies (e.g., Tully & Fouque 1985; Tully *et al.* 1998; Masters *et al.* 2003). We used the corrections suggested by Tully *et al.* (1998) and Masters *et al.* (2003), who obtained an empirical relation between the total luminosity and the dust content in disk galaxies. The extinction, A_λ^i , is given as a function of the axial ratio (b/a) in the pass-band λ as:

$$A_\lambda^i = \gamma_\lambda \log(a/b), \quad (3)$$

where a and b are the length of the semi-major and the semi-minor axis respectively. The γ was estimated in each band using an equation of the following form:

$$\gamma_\lambda = c_0 + c_1(M_\lambda + c_2), \quad (4)$$

where the values of the coefficients c_i ($i = 0, 1, 2$) for each band and for different magnitude ranges are given in Table 2. M_λ is the absolute magnitude in the corresponding pass-band.

5. Sample selection and the error budget

The slope and the scatter in the TF relation depends on the selection criteria and biases in the sample. Bernstein *et al.* (1994) outlined some criteria for galaxies to be included in TF studies. They preferred non-interacting galaxies of types Sb–Sd with steep HI profiles at the end velocities, having smooth outer isophotes and without a prominent bar. Galaxies with inclination $< 45^\circ$ are generally not included in TF studies. This is

Table 2. Coefficients for internal dust extinction correction.

Band	mag. range	c_0	c_1	c_2
R	> -16.2	0.00	0.00	0.00
	< -16.2	0.00	-0.24	16.2
J	> -22.2	0.77	-0.26	23.0
	< -22.2	0.60	-0.04	23.0
H	> -22.2	0.58	-0.26	23.0
	< -22.2	0.38	-0.01	23.0
K	> -22.2	0.31	-0.13	23.0
	< -22.2	0.20	0.00	23.0

because for low inclination galaxies, it is difficult to get an accurate estimate of the position angle and the inclination angle. This results in uncertain deprojected rotation velocities. For instance, for a TF slope of -10 in the near-infrared, a marginal error of 5° in the inclination angle introduces an additional scatter of nearly 0.4 mag, 0.2 mag and 0.05 mag in the TF relations at inclinations of 45° , 60° , and 80° respectively. At the same time, nearly edge-on galaxies will have their optical luminosities poorly estimated due to uncertainties in the internal dust extinction corrections. Therefore, TF studies on galaxies having intermediate inclination angles, and in the near-infrared bands are preferred. One would also like to reduce scatter due to uncertainties in the distances to galaxies in a sample. For this reason, galaxies in a group or a cluster may be the right choice instead of field galaxies.

The survey of the Eridanus group was not volume-limited. The rotation curves were measured for about half of the spiral galaxies. Therefore, somewhat relaxed criteria compared to those described above are used to construct the sample for the present study. The S0/a and interacting galaxies were excluded from the sample. Galaxies with inclination less than 35° are not included. There are 4 galaxies (ESO 548-G 065, ESO 549-G 002, IC 1962, NGC 1422), whose rotation curves are nearly flat. The sample consists of finally a total of 17 galaxies for which the HI rotation curves are flat or nearly flat in the outer regions. These galaxies are included in the sample. The photometric data on all galaxies are not available in all the bands. The photometric and the kinematical properties of the galaxies in the Eridanus group are given in Table 3.

We have assumed all the galaxies to be at a distance of $\sim 23 \pm 2$ Mpc. This assumption is based upon the distance estimates using the surface brightness fluctuations in the I-band and in the K-band of a few early type galaxies in the different sub-groups of the Eridanus group (Jensen *et al.* 1998; Tonry *et al.* 1997, 2001). The mean distance modulus ($m - M$) of NGC 1400 (S0, 558 km s^{-1}), NGC 1407 (E, 1779 km s^{-1}), NGC 1395 (E, 1717 km s^{-1}), NGC 1332 (S0, 1524 km s^{-1}), and NGC 1426 (E, 1443 km s^{-1}) is 31.8 ± 0.2 mag.

The photometric precision is estimated taking into account the measurement uncertainties. For typical magnitudes of the Eridanus galaxies ($m \sim 12$ mag) in the R band, the photometric accuracy is estimated as ~ 0.2 mag (Omar 2004). The accuracy of the 2MASS photometry is taken as 0.1 mag. There seems to be not much understanding on how much additional scatter can arise in the TF relations due to uncertainties in the internal dust extinction. Verheijen (2001) noted that the scatter in the TF relation

Table 3. Sample of galaxies used to construct the TF relation.

Name	Type	Incl.	$\log(W)$	M_R	M_J	M_H	M_K
ESO 482-G 005	SBdm	84.6	2.205	–	–17.31	–	–
ESO 482-G 035	SBab	51.8	2.374	–18.86	–20.62	–21.08	–21.45
ESO 548-G 021	SBdm	88	2.244	–	–18.47	–18.19	–18.99
ESO 548-G 065	SBa	79.6	2.164	–16.96	–18.29	–	–
ESO 549-G 002	IBrm	53.1	2.048	–17.46	–18.33	–	–
IC 1952	SBbc	81.4	2.427	–19.39	–21.55	–22.09	–22.19
IC 1953	SBd	50.6	2.476	–20.18	–22.13	–22.68	–23.04
IC 1962	SBdm	79.8	2.216	–17.64	–18.24	–18.21	–19.41
MCG -03-10-041	SBdm	62.5	2.342	–18.25	–19.92	–20.44	–20.54
NGC 1325	SAbc	74.5	2.500	–20.69	–22.64	–23.18	–23.33
NGC 1371	SABa	48.1	2.716	–21.12	–23.38	–24.00	–24.21
NGC 1385	SBcd	51.5	2.442	–20.97	–22.80	–23.41	–23.55
NGC 1414	SBbc	79.8	2.193	–16.95	–19.27	–19.68	–20.10
NGC 1422	SBab	79.6	2.169	–17.95	–20.38	–20.95	–21.16
SGC 0321.2-1929	IBrm	36.2	2.052	–15.04	–	–	–
UGCA 068	SABcdm	37.6	2.244	–17.57	–18.90	–19.15	–19.65
UGCA 077	SBdm	64.3	2.193	–17.30	–	–	–

Notes: **Column 1:** Name of the galaxy. **Column 2:** Hubble type. **Column 3:** The inclination angle (in degree) as obtained from the tilted ring model or obtained from the R-band optical images in those cases where the tilted ring model did not provide a satisfactory fit. **Column 4:** Width W is twice the value of the velocity at the flat part of the rotation curve or, of the rotation velocity at the last measured point. **Columns 5–8:** Absolute total magnitudes in the R, J, H, and K bands respectively. These magnitudes include all the corrections as described in section 4. All the galaxies are assumed to be at a distance of 23 Mpc (distance modulus ~ -31.8).

is increased by a marginal amount (~ 0.04 mag) from the K band to the R band. If we attribute this extra scatter to the extinction corrections, it appears that the internal dust extinction corrections are fairly well determined in these bands. There may be a significant contribution to the total scatter due to errors in the inclination angles of galaxies as explained before. We estimate the error budget for an intermediate inclination of 60° . Though all the errors may not be truly random in nature, we estimate the total expected rms assuming that the different errors add up randomly. This combined error is estimated as 0.3 mag.

6. Tully–Fisher relations

6.1 Classical luminosity–linewidth relation

Figure 2 shows TF relations in the R, J, H, and K bands. A least square fit is carried out to obtain the slope ‘ a ’ and the intercept ‘ b ’ for an equation of the form $\text{mag} = a \log(W) + b$. The fit is carried out using data for galaxies having flat rotation curves ($W = 2V_{\text{flat}}$). All such data points are given equal weights. The total observed scatter (rms of residuals) is estimated in magnitude. The slope, the intercept, and the scatter obtained from the fit are given in Table 4.

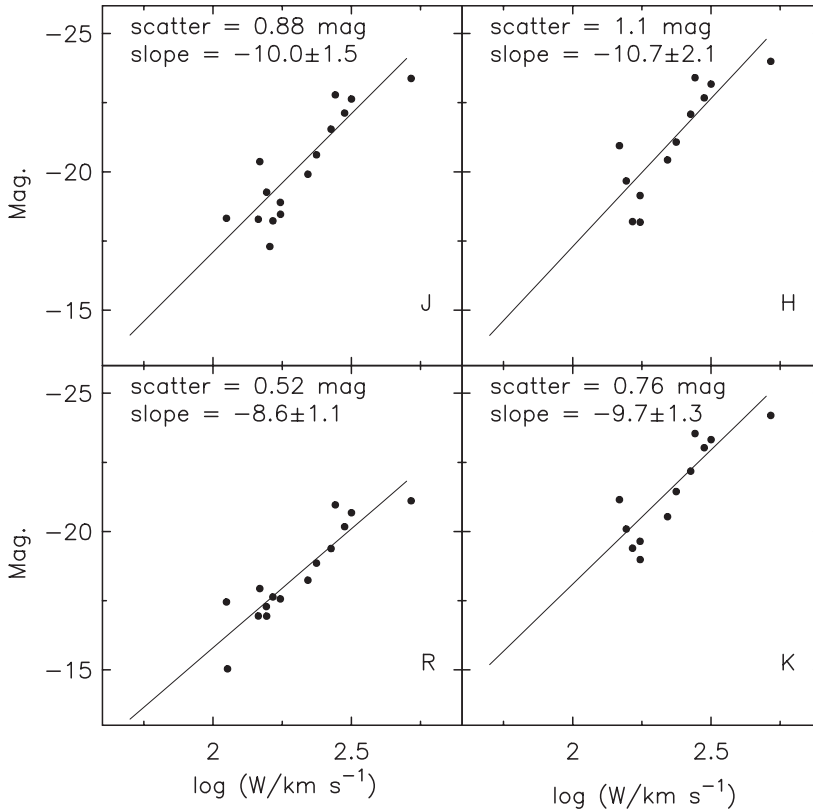


Figure 2. The Tully–Fisher relations in the R, J, H, and K bands. The values of $\log W$ (abscissa) and absolute magnitude (ordinate) are from Table 3.

Table 4. The TF parameters.

Band	Slope	Intercept	Scatter
Classical TF			
R	-8.6 ± 1.1	1.4 ± 2.5	0.52 mag
J	-10.0 ± 1.5	2.9 ± 3.5	0.88 mag
H	-10.7 ± 2.1	4.1 ± 5.0	1.10 mag
K	-9.7 ± 1.3	1.3 ± 4.1	0.76 mag
Baryonic TF			
R	4.1 ± 0.7	0.4 ± 1.6	0.19
J	4.1 ± 0.7	0.4 ± 1.6	0.18
H	3.7 ± 0.8	1.5 ± 2.0	0.18
K	3.5 ± 0.8	1.8 ± 1.8	0.16

Note: The scatter is the reduced chi-square value of the fit.

The slopes of the TF relations obtained in the present analysis are comparable to those obtained in other studies of nearby groups and clusters (e.g., Giovanelli *et al.* 1997; Verheijen 2001). The TF slopes from the current study are -8.6 ± 1.1 , -10.0 ± 1.5 , -10.7 ± 2.1 , -9.7 ± 1.3 in the R, J, H, and K bands respectively. The slopes obtained by Verheijen (2001) for the Ursa–Major galaxies are -8.3 ± 0.6 and -11.2 ± 0.7 in the R and K bands respectively. They reported a scatter as low as 0.21 mag for the K-band. Giovanelli *et al.* (1997) obtained a mean slope of -7.6 in the I band for a set of nearby groups and clusters. They also reported an I band TF slope of -7.88 ± 0.56 with a scatter of 0.36 mag for the Eridanus group using their single dish HI data and the optical rotation curves. They pointed out that both the slope and the scatter for the Eridanus group are somewhat different from the average values for other groups and clusters in their sample. Recently, using the photometric data from 2MASS and the kinematical data from Giovanelli *et al.* (1997), Bamford (2002) reported slopes of -10.8 , -10.9 , and -11.1 for the J, H, and K bands respectively, and a scatter of 0.57 mag for the Eridanus galaxies. It can be seen from Table 4 that the scatter in the TF relations in the present study (0.5–1.1 mag) is also significantly higher compared to that obtained for the Ursa–Major group and for other nearby groups and clusters.

The larger scatter in the TF relations for the Eridanus galaxies in various studies including those based on the near-infrared data indicate that there are variations in the distances to the galaxies in the group. If the additional scatter in the TF relation is attributed to the depth of the group then it indicates that the depth of the Eridanus group can be as large as 4 Mpc. Perhaps, it is tied up with the dynamical stage of the Eridanus group where galaxies or small groups are being accreted from outer regions (Willmer *et al.* 1989).

6.2 Baryonic TF relation

The baryonic TF relations are constructed using the combined HI and Helium gas mass, and the stellar mass obtained using the stellar luminosities in different bands and their corresponding M^*/L ratios from the spectrophotometric population synthesis models of Bell & de Jong (2001). These models predict a linear relationship between the colours of galaxies and the M^*/L ratio for a band. These models allow the estimation of

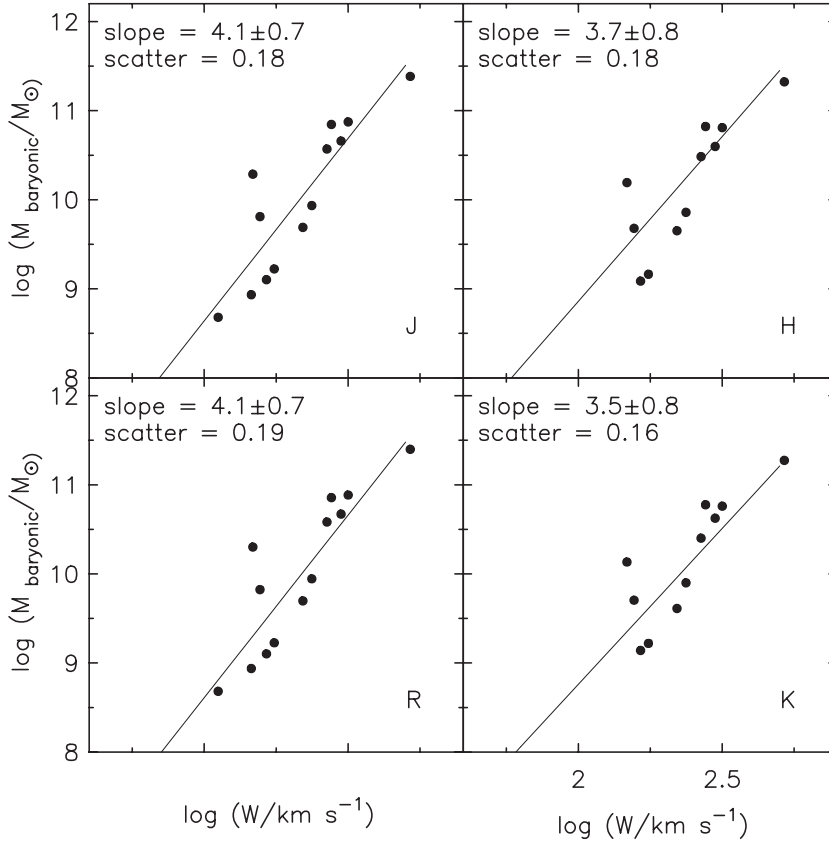


Figure 3. The baryonic TF relations constructed using the combined HI and Helium gas mass, and the stellar mass obtained using the stellar luminosities in the R, J, H, and K bands and their corresponding M^*/L ratios from the models of Bell & de Jong (2001).

the total stellar mass from the stellar luminosity in any wave-band provided the colours of galaxies are known. For the current study, the total stellar masses are estimated using the (R–J) colours. The Helium mass is taken as 1.32 times the HI mass. Since the molecular and the ionized gas masses are unknown, the baryonic masses thus obtained are lower limits. The baryonic TF relations are plotted in Fig. 3.

If the TF relations were truly a correlation between the baryonic mass and the rotation velocity, and the models of Bell & de Jong (2001) make accurate predictions of M^*/L for each band, the baryonic TF slope should be independent of pass-band. A nearly pass-band independent baryonic TF relation is obtained in the R, J, H, and K bands as indicated in Fig. 3 and Table 4. The slope of the Baryonic TF relation is in the range 3.5–4.1.

McGaugh *et al.* (2000) obtained a slope of 3.98 ± 0.12 for a sample of galaxies with rotation velocities over a large range ($30\text{--}300 \text{ km s}^{-1}$). Using data from Verheijen (2001), Bell & de Jong (2001) obtained a slope of 3.5 ± 0.4 for the Ursa–Major group of galaxies. The slope of the baryonic TF relation is predicted to be 3 for non-baryonic dark matter (Navarro & Steinmetz 2000) and 3.5 for the collisional baryonic dark matter (Walker 1999). It is worth mentioning here that all these observational baryonic

TF relation (including the present one) do not include the molecular and the ionized gas masses. These mass components will be significant for dwarfs and gas rich low surface brightness (LSB) galaxies, whose rotation velocities are lower compared (typically $< 100 \text{ km s}^{-1}$) to that for high surface brightness (HSB) galaxies. The inclusion of the molecular and the ionized gas masses to the total baryonic mass in a sample having all types (dwarfs, LSB and HSB) of galaxies will make the slope of the baryonic TF relations shallow. Further, the M^*/L ratios in the models of Bell & de Jong (2001) were calibrated assuming that the galactic disks are supported mainly by rotation. If this is not the case, the baryonic TF relation will be even shallower. We believe that with uncertainties in these analyses and the incomplete information about the total baryonic mass in galaxies, currently no dark matter model can be ruled out based on the baryonic TF relation.

7. Conclusions

- The TF slopes in the Eridanus galaxies are consistent with those obtained in other studies of nearby groups and clusters. The slopes of the TF relations (absolute magnitude vs $\log 2V_{\text{flat}}$) for the galaxies in the Eridanus group are -8.6 ± 1.1 , -10.0 ± 1.5 , -10.7 ± 2.1 , and -9.7 ± 1.3 in the R, J, H, and K bands respectively.
- The TF relation for the Eridanus group has a significantly larger scatter ($\sigma \sim 0.5 - 1.1 \text{ mag}$) compared to that for other groups and clusters ($\sigma \sim 0.2 - 0.4 \text{ mag}$).
- If the larger scatter in the TF relations indicates variations in the distances to galaxies in a sample, the maximum difference in the distances to galaxies in the Eridanus group will be $\sim 4 \text{ Mpc}$. This implies that the galaxies in the Eridanus group are more dispersed in space compared to other groups and clusters. It is suspected that we are looking down a filament.
- The value of the index α ($M_{\text{baryon}} \sim V_{\text{flat}}^\alpha$) of the baryonic TF relation is in the range 3.5–4.1.

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Appendix follows

Appendix 1

Table 5. Photometric results.

Galaxy	Type P.A.± Incl.± r(B ₂₅) (kpc)	F	r _d ± (kpc)	r ₍₂₅₎ r ₍₂₁₎ r _(20.5) r ₍₂₀₎ (kpc)	μ ⁰ μ ^{0,i} ± (mag arcsec ⁻²)	m ₂₆ m ₂₁ m _{20.5} m ₂₀ (mag)	m ^{total} (mag)
E 482-005	SBdm	R	—	—	—	—	14.50
	79 0	J	1.89 0.10	2.5	19.51 22.06 0.05	15.10	—
	85 2	H	—	—	—	—	—
	5.6	K	—	—	—	—	—
E 482-035	SBab	R	2.20 0.03	7.4	21.02 21.23 0.02	13.19	12.94
	4 1	J	1.95 0.05	3.6	18.89 19.39 0.04	11.76	11.18
	52 1	H	1.61 0.08	3.6	18.13 18.64 0.08	11.18	10.72
	4.4	K	1.96 0.12	3.4	18.14 18.65 0.09	11.03	10.35
E 548-021	SBdm	R	—	—	—	—	—
	70 1	J	4.71 0.20	5.6	19.40 23.01 0.02	14.12	13.33
	88 0	H	3.37 0.23	4.5	18.86 22.48 0.05	14.15	13.62
	6.7	K	3.71 0.34	4.5	18.43 22.06 0.06	13.68	12.81
E 548-065	SBa	R	1.77 0.01	5.8	21.34 22.11 0.01	15.31	14.85
	41 0	J	1.74 0.10	2.2	18.99 20.78 0.05	13.98	13.51
	80 1	H	—	—	—	—	—
	5.0	K	—	—	—	—	—
E 549-002	IBm	R	1.51 0.03	4.5	21.79 21.88 0.03	14.73	14.34
	32 1	J	1.08 0.15	1.3	19.42 19.91 0.21	14.28	13.48
	53 1	H	—	—	—	—	—
	4.4	K	—	—	—	—	—
I 1952	SBbc	R	3.52 0.06	11.2	20.21 21.21 0.02	13.05	12.42
	140 0	J	3.60 0.04	9.0	17.88 19.91 0.01	10.78	10.26
	81 0	H	3.46 0.06	9.0	17.17 19.22 0.02	10.15	9.72
	8.7	K	3.49 0.07	7.8	16.92 18.97 0.02	9.98	9.62

I 1953	SBd	R	4.20	0.11	12.1	21.08	21.26	0.03	11.95	11.62
	121 0	J	3.78	0.10	7.4	19.03	19.50	0.03	10.39	9.68
	51 1	H	3.78	0.07	7.6	18.42	18.90	0.02	9.75	9.13
	9.4	K	4.19	0.13	7.2	18.22	18.70	0.03	9.51	8.77
I 1962	SBdm	R	2.77	0.05	9.2	21.64	22.54	0.02	14.62	14.16
	1 1	J	2.06	0.14	2.5	19.70	21.56	0.07	14.45	13.57
	80 2	H	1.76	0.21	2.0	19.29	21.16	0.14	14.44	13.60
	9.1	K	1.68	0.18	2.7	18.32	20.19	0.13	13.19	12.40
M-03-10-041	SBdm	R	2.67	0.08	6.9	21.81	22.09	0.04	13.99	13.56
	166 1	J	3.13	0.28	2.0	20.14	20.93	0.10	13.54	11.89
	63 1	H	2.68	0.26	2.7	19.36	20.17	0.12	12.59	11.37
	6.7	K	2.36	0.29	2.0	18.95	19.77	0.16	12.63	11.26
N 1325	SAbc	R	4.93	0.03	19.3	20.45	21.14	0.01	11.75	11.12
	53 1	J	4.60	0.06	11.7	18.04	19.46	0.01	9.74	9.17
	75 0	H	4.51	0.09	12.3	17.36	18.78	0.03	9.06	8.63
	15.8	K	4.61	0.09	10.8	17.13	18.56	0.02	8.88	8.48
N 1371	SABa	R	3.55	0.02	17.2	20.22	20.38	0.01	10.98	10.69
	134 1	J	3.34	0.02	10.3	17.65	18.07	0.01	8.75	8.42
	48 1	H	3.28	0.02	10.8	16.96	17.38	0.01	8.06	7.81
	18.8	K	3.37	0.04	9.6	16.76	17.19	0.01	7.87	7.60
N 1385	SBcd	R	3.32	0.04	15.2	20.20	20.42	0.02	11.14	10.84
	173 1	J	3.21	0.10	8.7	18.10	18.60	0.05	9.43	9.02
	51 1	H	3.16	0.10	9.2	17.39	17.89	0.05	8.72	8.40
	11.4	K	2.98	0.11	8.1	17.13	17.64	0.07	8.60	8.26
N 1414	SBbc	R	1.88	0.04	5.6	21.16	22.07	0.02	15.18	14.86
	170 1	J	1.96	0.06	3.8	18.69	20.55	0.03	12.95	12.54
	80 1	H	1.57	0.04	4.0	17.85	19.72	0.04	12.37	12.13
	5.8	K	1.81	0.07	3.4	17.73	19.60	0.04	12.23	11.71
N 1422	SBab	R	2.32	0.03	9.2	20.90	21.80	0.02	14.30	13.86
	66 0	J	2.44	0.03	5.8	18.29	20.12	0.01	11.87	11.43
	80 1	H	2.12	0.05	5.8	17.38	19.22	0.03	11.15	10.85
	7.4	K	1.84	0.03	5.4	16.94	18.79	0.02	10.98	10.65

Table 5. (Continued)

Galaxy	Type P.A. $\pm$ Incl. $\pm$ r(B ₂₅) (kpc)	F	r _d $\pm$ (kpc)	r ₍₂₅₎ r ₍₂₁₎ r _(20.5) r ₍₂₀₎ (kpc)	μ^0 (mag arcsec ⁻²)	$\mu^{0,i} \pm$ (mag arcsec ⁻²)	m ₂₆ m ₂₁ m _{20.5} m ₂₀ (mag)	m ^{total} (mag)		
S 0321.2-1929	IBm	R	2.10	0.20	0.2	24.79	24.83	0.065	17.99	16.76
	170 0	J	—	—	—	—	—	—	—	—
	36 6	H	—	—	—	—	—	—	—	—
	4.4	K	—	—	—	—	—	—	—	—
U 068	SABcdm	R	2.06	0.05	5.2	22.41	22.45	0.03	14.53	14.24
	42 1	J	1.29	0.12	1.3	19.58	19.80	0.13	13.84	12.90
	38 3	H	1.43	0.22	1.3	19.27	19.51	0.20	13.75	12.66
	5.7	K	1.08	0.16	1.3	18.55	18.79	0.24	13.08	12.15
U 077	SBdm	R	2.42	0.05	6.0	22.52	22.83	0.02	14.97	14.51
	138 1	J	—	—	—	—	—	—	—	—
	64 1	H	—	—	—	—	—	—	—	—
	5.7	K	—	—	—	—	—	—	—	—

Notes: **Column 1:** The name of the galaxy in an abbreviated format. The ESO galaxies are written as E, NGC is written as N, UGCA is written as U, IC is written as I, MCG is written as M, and SGC is written as S. **Column 2:** The morphological type, position angle and inclination with associated errors in the estimation determined from the R band observations or from DSS images when R band observations were not available, radius in kpc at 25 mag arcsec⁻² B band isophote taken from RC3. The P.A. is measured from North to East (counter-clockwise). **Column 3:** The filter/band. **Columns 4 & 5:** The disk scale length and associated errors in the fitting procedure in units of kpc. **Column 6:** The radius in kpc corresponding to the isophotal level of 25 mag arcsec⁻² in the R, 21 mag arcsec⁻² in the J, 20.5 mag arcsec⁻² in the H, and 20 mag arcsec⁻² in the K band. **Columns 7, 8 & 9:** The central surface brightness of the disk as obtained from the fit, and the values corrected for the inclination, Galactic extinction, and internal extinction. The error is obtained from the fitting procedure and does not include calibration errors. **Column 10:** The isophotal magnitude of the galaxy estimated within the fixed aperture at 26.0 mag arcsec⁻² for the R band, 21.0 mag arcsec⁻² in the J band, 20.5 mag arcsec⁻² in the H band, and 20.0 mag arcsec⁻² in the K band. **Column 11:** The total magnitude extrapolated to infinity with corrections for Galactic and internal extinctions.

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Variable Quasi Periodic Oscillations during an Outburst of the Transient X-ray Pulsar XTE J1858 + 034

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Abstract. We have investigated the Quasi Periodic Oscillation (QPO) properties of the transient accreting X-ray pulsar XTE J1858 + 034 during the second outburst of this source in April–May 2004. We have used observations made with the Proportional Counter Array (PCA) of the Rossi X-ray Timing Explorer (RXTE) during May 14–18, 2004, in the declining phase of the outburst. We detected the presence of low frequency QPOs in the frequency range of 140–185 mHz in all the RXTE-PCA observations. We report evolution of the QPO parameters with the time, X-ray flux, and X-ray photon energy. Though a correlation between the QPO centroid frequency and the instantaneous X-ray flux is not very clear from the data, we point out that the QPO frequency and the one day averaged X-ray flux decreased with time during these observations. We have obtained a clear energy dependence of the RMS variation in the QPOs, increasing from about 3% at 3 keV to 6% at 25 keV. The X-ray pulse profile is a single peaked sinusoidal, with pulse fraction increasing from 20% at 3 keV to 45% at 30 keV. We found that, similar to the previous outburst, the energy spectrum is well fitted with a model consisting of a cut-off power law along with an iron emission line.

Key words. Stars: pulsar: individual (XTE J1858 + 034)—X-rays: pulsar.

1. Introduction

Among the accretion powered X-ray pulsars, Quasi Periodic Oscillations (QPO) were first discovered in the X-ray flux of the transient source EXO 2030 + 375 (Angelini *et al.* 1989). Since then, QPOs have been detected in the X-ray power density spectra of almost a dozen accretion-powered pulsars (Psaltis 2004). QPO centre frequencies as seen in the pulsars vary widely, ranging from 1 mHz to 40 Hz. In the models that deal with QPOs in X-ray pulsars, a correlation is expected between the QPO centre and the X-ray flux, which in turn is determined by the mass accretion rate. Hence, intensity dependent variable QPOs can provide an interesting insight into the physics of accretion. However, QPOs are in general rare and transient events in X-ray pulsars.

To date, six out of the dozen X-ray Binary Pulsars (XBP) showing QPOs are transient sources (Psaltis 2004). XTE J1858 + 034 is one of the transient XBPs in which QPOs have been detected. This X-ray binary was discovered with the RXTE All Sky Monitor (ASM) in 1998 February during a transient outburst (Remillard & Levine 1998). Observations made with the Proportional Counter Array (PCA) of the RXTE during the first outburst of the source showed single peaked strong sinusoidal pulsations with a period of ~ 221.0 s (Takeshima *et al.* 1998; Paul & Rao 1998). QPOs at 110 mHz with about 7% RMS fluctuation were discovered from the same observation (Paul & Rao 1998). The energy spectrum was measured to be hard, similar to the spectra of XBPs. From the transient nature of this source, its hard X-ray spectrum, and presence of pulsations, it was suggested to be a Be-X-ray binary (Takeshima *et al.* 1998) which is also supported from the optical observations of H_α emission line (Reig *et al.* 2004).

The source had a second outburst that was first detected with INTEGRAL in March 2004 and subsequently again in April 2004 (Molkov *et al.* 2004). The RXTE-ASM also detected the outburst and it was followed up with the RXTE-PCA. We detected the QPO features again during this second outburst from an observation made on May 14, 2004. Subsequently, a series of target of opportunity observations of XTE J1858 + 034 were made with RXTE during May 16–18 to study the evolution of the QPO properties. We have investigated the evolution of the QPO parameters with time, X-ray flux and X-ray photon energy. X-ray pulsation properties and the X-ray spectrum were also measured from these observations. In the following sections, we present the X-ray observations, analyses and results and a discussion based on the results obtained.

2. Observations

The observations of XTE J1858 + 034 reported here were made during May 14–18, 2004 with the pointed mode instruments of the RXTE satellite. The RXTE-ASM lightcurve of XTE J1858 + 034 is shown in Fig. 1 for the period of March–May 2004 with a bin-size of one-and-a-half days. The ASM light curve clearly shows that during the RXTE pointed observations, the source was in the decaying phase of this outburst.

The RXTE satellite carries a large area Proportional Counter Array (PCA), consisting of five Xenon filled proportional counter units (PCUs) sensitive in the energy range of 2–60 keV with a total effective area of 6250 cm² (Jahoda *et al.* 1996). It has also a set of high energy crystal scintillation experiment (HEXTE; 15–200 keV; 1600 cm² area), and a continuously scanning all-sky monitor (ASM; 2–10 keV; 90 cm²). A full description of the ASM detector can be found in Bradt *et al.* (1993). Here we have used the RXTE-ASM long term light curve of XTE J1858 + 034 available from the beginning of the RXTE mission in 1995 to the present and data from six pointed observations with RXTE-PCA. The details of the observations used in the present work have been given in a nutshell in Table 1. In total, there were data from eight orbits of the satellite. Two of the six observations, Obs. B & F had two orbits of data each (B1, B2 & F1, F2 in Table 1). The total good time interval (GTI) with the PCA from these six observations was 18.7 ks. Figure 1 shows the RXTE-ASM long term light curve (bin size of one-and-a-half days for clarity) along with the background subtracted RXTE-PCA count rates superposed on it. Though the beginning of the outburst is not clear, the gradual fall in the long term intensity can be clearly seen from the figure.

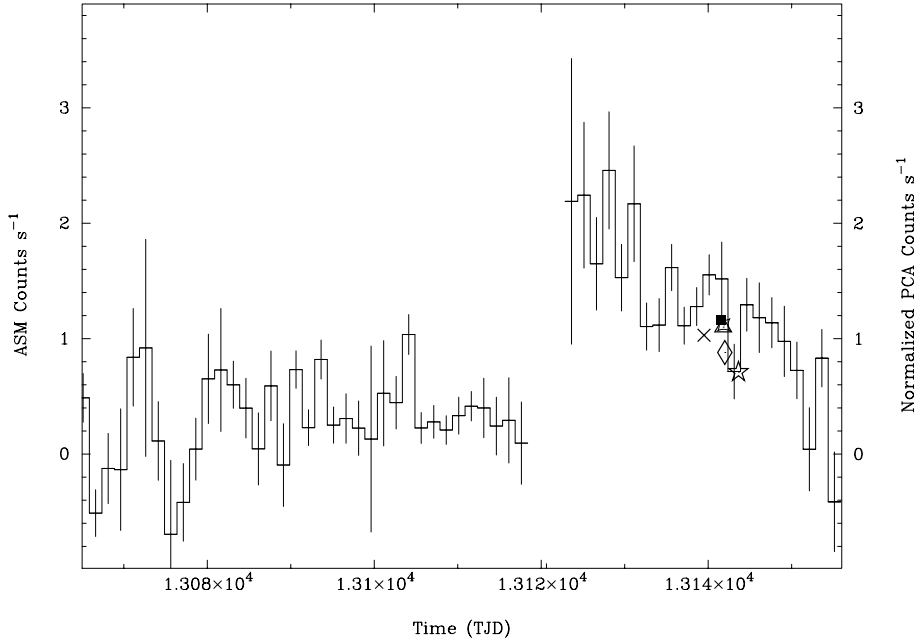


Figure 1. The section of the RXTE-ASM long term light curve depicted between March 1 and May 31, 2005, showing the 2004 outburst in detail. The bin size is one-and-a-half days. The six PCA observations are shown (with different markers) with their respective background subtracted count rates (normalized by the mean count rate).

Table 1. Observation log.

Observations	Date	Start time (UT)	Stop time (UT)	Exposure (s)	Count rate (s ⁻¹)
A	2004-05-14	11:30	12:26	3344	365
B1	2004-05-16	11:14	11:42	1665	422
B2	2004-05-16	13:54	14:48	3213	428
C	2004-05-16	18:47	19:21	2017	399
D	2004-05-16	20:24	20:45	1225	408
E	2004-05-16	23:41	00:05	1403	331
F1	2004-05-18	13:11	14:06	3297	279
F2	2004-05-18	14:43	15:25	2517	277

3. Data analysis and results

3.1 Light curve and pulse profile

From all the eight segments in the six observations, we extracted the light curves, each with a bin size of 0.125 s and then applied Barycentre correction on them. All the individual light curves showed the pulses. The combined background subtracted light curve of all the six observations binned at the pulse period of ~ 220 s had an average count rate of ~ 260 counts s⁻¹ and also had a gradual fall in intensity with time (see

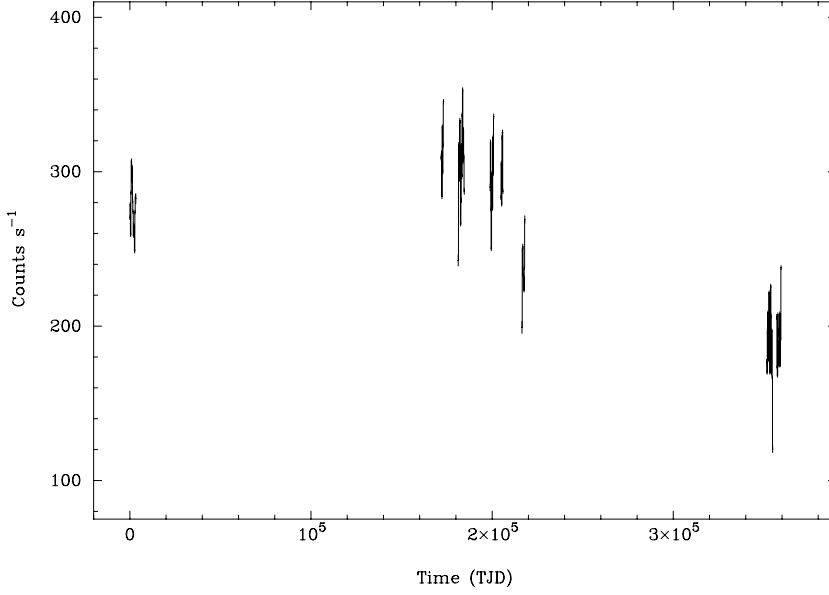


Figure 2. The RXTE-PCA background subtracted light curve in the energy band of 2–60 keV obtained from the 2004 observations is shown here with a bin size of 220 s.

Fig. 2). The pulse folding and the χ^2 maximization method gave a pulse period of 219.823 ± 0.005 s within 90% confidence interval. A slow transient pulsar is expected to have a large spin-up rate in a high intensity state like this. However, a small duty cycle during the five days of observation does not allow us to measure the spin-up rate. The 2–60 keV X-ray light curve, obtained from the PCA detectors is shown in Fig. 2. This figure is generated with a bin size of 220 s, same as the spin period, which removes the pulsation related intensity variations. Though the overall intensity changed with time, there was no significant intensity variation at a time scale of a few thousand seconds.

We have folded the energy resolved background subtracted light curves with the pulse period mentioned above and the resultant pulse profiles are shown in Fig. 3 for observation A. The pulse profiles shown for six out of the eight energy bands are single peaked sinusoidal in energy bands up to 25–35 keV range with a pulse fraction (defined as Maximum–Minimum/Maximum) gradually increasing from 20% at 3 keV to 45% at 30 keV. From Fig. 3, we see that pulsations exist up to 45 keV also with a pulse fraction of 35% in the energy band of 35–45 keV. Pulsations above 45 keV is not detectable with the PCA due to low signal-to-noise ratio and low detection efficiency.

3.2 Power density spectrum and related results

We have generated the Power Density Spectra (PDS) from the 2–60 keV 0.125 s time resolved light curves from each orbit of the satellite. The light curves were broken into segments of length 128 s and the PDS obtained from each of these segments were averaged to produce the final PDS for data from each orbit of the satellite. A broad QPO feature between 100 mHz and 200 mHz was very prominent in all the PDS. The

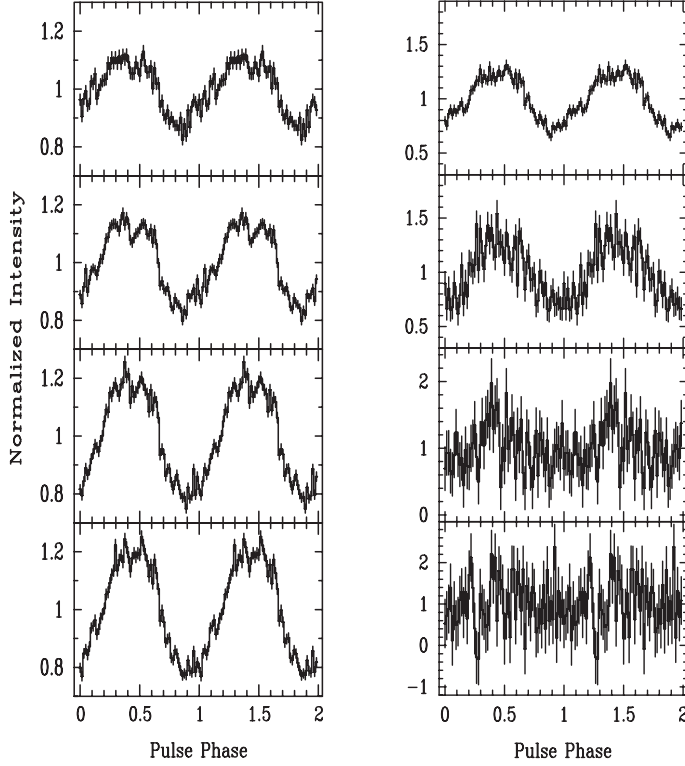


Figure 3. The nearly sinusoidal pulse profiles of XTE J1858 + 034 folded at a period of 219.8 s are shown (left, from the top) for 2–5 keV (first panel), 5–8 keV (second panel), 8–12 keV (third panel) and 12–18 keV (fourth panel). In the right the same is shown for 18–25 keV (top panel), 25–35 keV (second panel), 35–45 keV (third panel) and 45–60 keV (bottom panel) energy bands. The bottom-right panel shows that pulsations are not detectable with the PCA above 45 keV.

continuum of the PDS in the frequency range of 7 mHz to 2 Hz fits well with a model consisting of a power law. The prominent QPO feature is fitted well with a Gaussian model component. In most of the PDS, there is also an evidence for a second QPO peak, probably a harmonic of the fundamental. However, we have refrained from quantifying the second peak as it has very low statistics. In Fig. 4, we have shown the power density spectrum from observation C which has a power law index of -0.93 and the Gaussian, QPO feature centered at 160 mHz. Evolution of the QPO frequency and the associated RMS for the eight data segments are shown in Fig. 5. While the QPO frequency decreased with time from 185 mHz on May 14 to about 140 mHz on May 18, the RMS variability associated with the QPO did not have any significant change. The RMS variability in the QPO feature is of the order of 4.0% to 3.0%. In order to investigate the energy dependence of the QPO feature, we generated PDS in the five energy bands of 2–5 keV, 5–8 keV, 8–12 keV, 12–18 keV and 18–30 keV respectively for Obs. A. We did not find any evidence of a QPO beyond the 18–30 keV band though. The PDS in the different energy bands were more or less identical in shape, and were fitted with the same model as described above. The RMS variation in the QPO feature in different X-ray energy bands obtained from observation A is

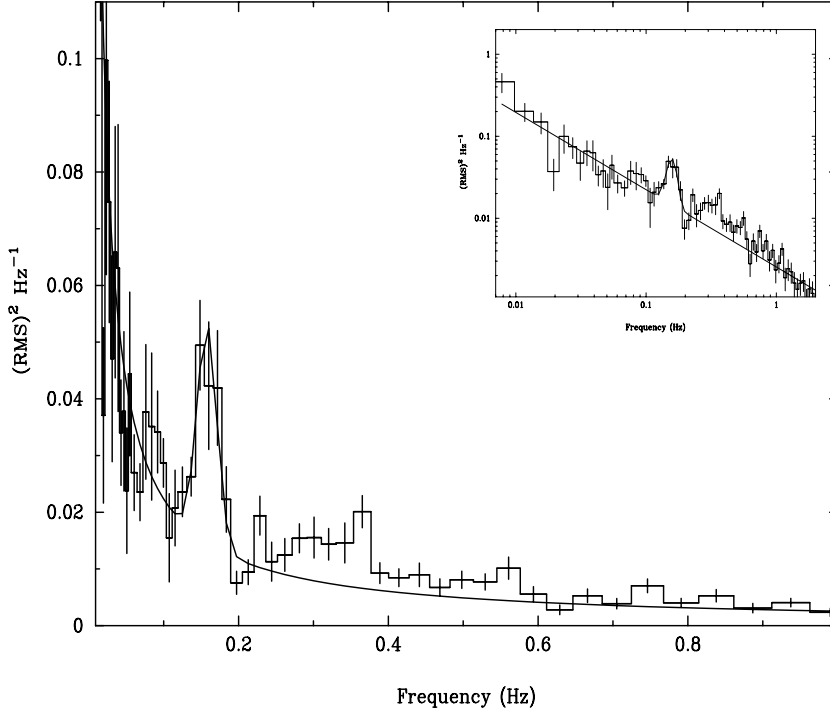


Figure 4. The power density spectrum of XTE J1858 + 034 generated from the 0.125 s binned light curve of observation C over the entire energy band of the PCA (in linear scale). The line represents the best fitted model in the frequency range of 0.007–2.0 Hz (shown for a restricted range to highlight the QPO feature) consisting of a power law and a Gaussian centred at the QPO frequency. The figure in the inset is a log–log plot of the same fit.

plotted in Fig. 6. We have obtained a clear energy dependence of the RMS variation in the QPOs, increasing from about 3% at 3 keV to 6% at 25 keV. The QPO centroid frequency is plotted against the 3–30 keV RXTE-PCA flux in Fig. 7, which does not show a clear correlation between the two.

3.3 Energy spectrum

For spectral analysis, we used the Standard-2 dataset from PCA which is stored with 16 s time resolution. We first extracted the source plus background spectra from all the eight segments of observation. The background spectra were modeled using the **pcabackest** package provided by the XTE guest observer facility (GOF). On the average three PCUs were ON, and data from all the PCUs were added together to produce the spectrum. Channels corresponding to energy less than 3 keV and greater than 30 keV were ignored because of low signal-to-noise ratio. Although the background simulation model that we have used takes care of the internal background, emission from an extended source like the Galactic ridge or the cosmic X-ray background are not included in it. From a detailed observation of the Galactic ridge obtained using the PCAs (Valinia & Marshall 1998), we have estimated that about 1.0% of the observed flux can be accounted for by the Galactic ridge emission and a very small fraction by

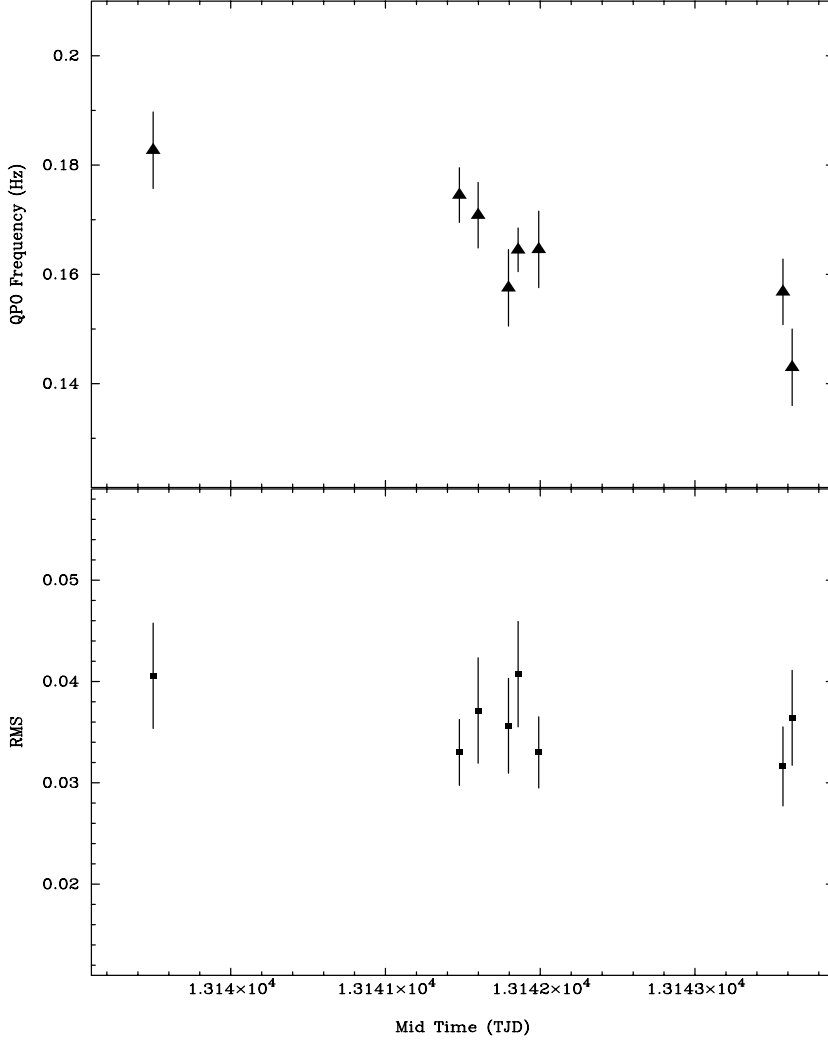


Figure 5. The QPO frequency and the RMS variation in the QPO feature are shown as function of time in the top and the bottom panels respectively.

the cosmic X-ray background. We have, therefore, explicitly incorporated the Galactic ridge emission as a sum of Raymond–Smith plasma and a power-law with appropriate normalizations. For fitting the pulse-phase averaged X-ray spectrum of the X-ray we have used a high energy cut-off power law along with a Gaussian emission line. The pulsar spectrum in the 3–30 keV range is found to be very hard with a photon index ~ 1.5 , cut-off energy of ~ 17 keV and e-folding energy of ~ 16 keV along with a neutral absorption with an equivalent Hydrogen column density of $\sim 10^{23}$ atoms cm^{-2} . The spectrum did not show any significant variation over the five days of observations. During these observations, the X-ray flux in the 3–30 keV range varied between 2.9×10^{-9} and 5.5×10^{-9} erg cm^{-2} s^{-1} . The X-ray spectrum from one of the observations is shown in Fig. 7 along with the best fit model and the residuals.

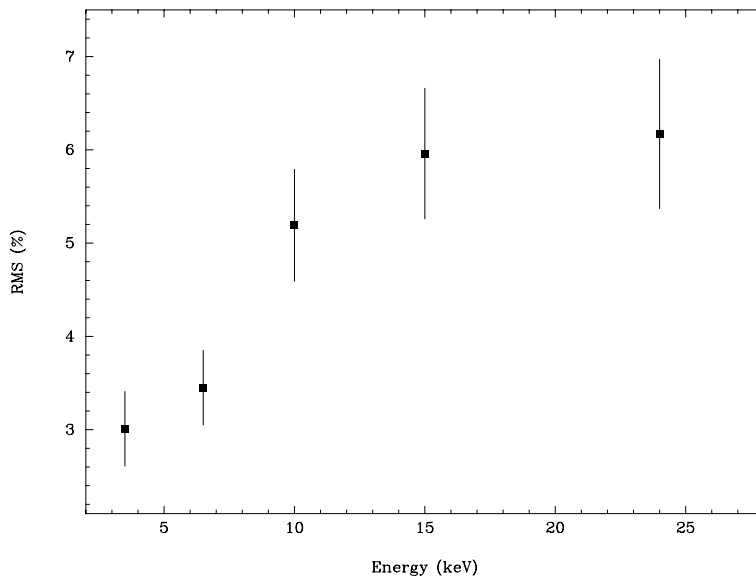


Figure 6. The RMS variation (in percentage) in the QPO feature determined from observation A is shown here as a function of the energy of the X-ray photons.

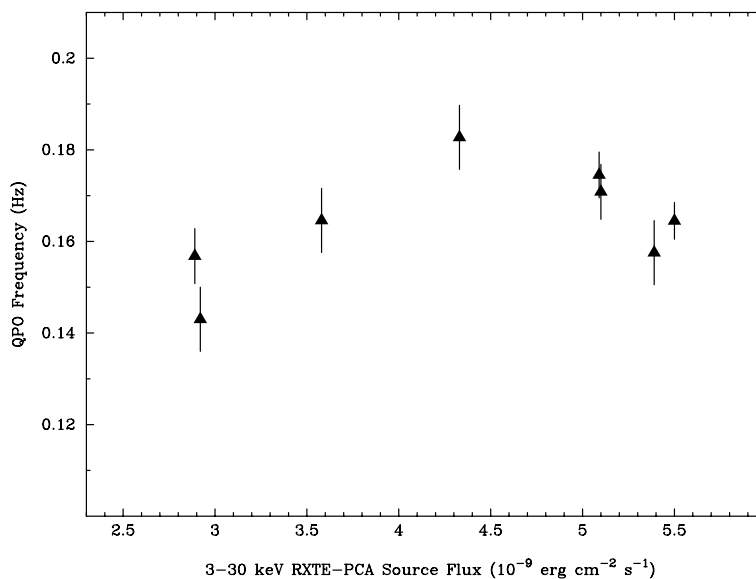


Figure 7. The QPO frequency measured from each segment of the RXTE-PCA light curve are shown here against the 3–30 keV X-ray flux.

4. Discussion

QPOs in accretion powered X-ray pulsars are in general explained with either the beat frequency model or the Keplerian frequency model. In the former (Alpar & Shaham 1985), blobs of matter orbit the Neutron Star (NS) at approximately the Keplerian

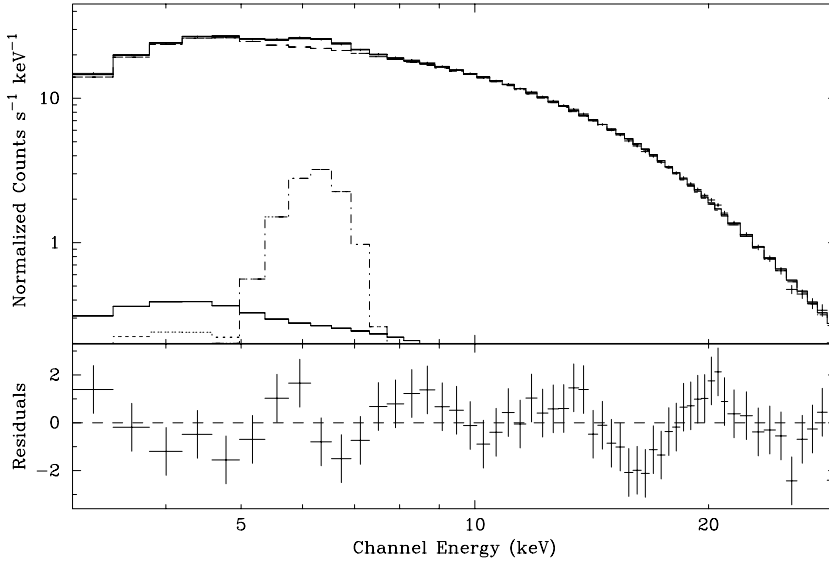


Figure 8. The X-ray spectrum obtained from observation B is shown here along with the best fit model components including a cut-off power law, a Gaussian emission line, and the additive background components as histograms. The residuals to the best fit model are shown in the lower panel in unit of their contribution to the total χ^2 .

orbital frequency of the inner edge of the accretion disk. The magnetic field rotates at the frequency of the NS and material inflow to the NS is modulated at the Keplerian frequency. If ν_k is the Keplerian frequency of the inner edge of the disk and ν_s be the spin frequency of the NS, then according to the beat frequency model, the frequency of the QPO (ν_{QPO}) is given by the relation $\nu_{QPO} = \nu_k - \nu_s$. According to the latter model (Van der Klis *et al.* 1987), the inner edge of the accretion disk contains structures that modulate the observed flux by obscuration and gives rise to a QPO at ν_k . A third model by Shirakawa & Lai (2002) is known as the magnetic disk precession model. It attributes the presence of QPOs in XBPs to the magnetically driven disk warping/precession, concentrated near the inner edge of the disk, at the magnetospheric boundary. The warping and precession of the disk arises from the magnetic torques which are present due to the interactions between the stellar field and the induced electric currents in the disk.

Out of the six transient XBPs showing QPOs, luminosity dependent QPO properties have been studied extensively for the two sources EXO 2030 + 375 (Angelini *et al.* 1989) and 3A 0535 + 262 (Finger *et al.* 1996) and both the beat and Keplerian frequency models are applicable to a good extent. In several XBPs with QPOs like Her X-1, Cen X-3, 4U 1626-67, LMC X-4 and SMC X-1, and two transients V 0332 + 52 and 4U 0115 + 63, the QPO frequency is lower than the pulsation frequency (Psaltis 2004). In these sources, the Keplerian frequency model is not applicable, since a Keplerian rotation frequency of the inner accretion disk that is smaller than the spin frequency of the neutron star will result into centrifugal inhibition of accretion. Again, for V 0332 + 52, the beat frequency model may also not be viable since the value of the magnetospheric boundary calculated from the QPO properties and from the observed luminosity do not agree (Takeshima *et al.* 1994). In the Low Mass XBP GRO

J1744-28, both these models are eliminated since in this source also $\nu_{QPO} < \nu_s$ and a flux change by a factor of ~ 7.5 did not have any corresponding change in QPO frequency.

The magnetic disk precession model (Shirakawa & Lai 2002) can explain the very low frequency QPOs in the XBPs, like in 4U 1626-67 (1 mHz) and may also explain the lowest frequency QPOs detected in some of the other XBPs (Her X-1, LMC X-4, and 4U 0115 + 63). While this model robustly accounts for QPOs with frequencies of the order of 1 mHz, the QPO frequency of XTE J1858 + 034 is two orders of magnitude higher and in this model, the production of QPOs at higher frequency depends on the details of the physics at the inner edge of the disk. Applicability of the magnetic disk precession model is therefore uncertain in this source.

It has already been mentioned that both the beat frequency and the Keplerian frequency models of QPOs predict a systematic increase of QPO frequency with the mass-accretion rate, which can be investigated by looking at the change of QPO frequency with the X-ray flux. This assumption is generally applicable to transient XBPs, in which the relation between X-ray luminosity and spin-up torque is well understood. The spin-up torque is in turn related to the mass accretion rate, inner radius of the accretion disk and rotation frequency of the inner disk. As the transient XBPs experience a wide range of mass-accretion rates, these are ideal objects to understand the QPO phenomena. In the case of XTE J1858 + 034 during the present outburst, we have detected changes in the QPO frequency over a span of five days. There is no discernible correlation between the QPO frequency and the instantaneous X-ray flux measured from the PCA data, at least during the observation duration (Fig. 7). However, as seen from the RXTE-ASM lightcurve, there is a gradual decrease in the overall one-and-a-half day averaged X-ray intensity and a gradual decrease in the QPO frequency (from 185 mHz to 140 mHz, Fig. 5) is also present during these observations. It may be possible that there is a correlation between the QPO frequency and the overall X-ray intensity or mass accretion rate which is not reflected in the short term. We note that in the RXTE-PCA observations of XTE J1858 + 034 made during the 1998 outburst, the source had a count rate (normalised for per detector unit) which was about a factor of 2.5–3.0 times less compared to the observations reported here. A smaller QPO frequency of 110 mHz during the 1998 outburst (Paul & Rao 1998) compared to the 140–185 mHz frequency during the 2004 outburst supports an inner accretion disk origin of the QPOs.

Absorption of X-rays by structures in the accretion disk should not be effective in producing hard X-ray QPOs and this process should definitely be more effective at lower X-ray energy. The detection of QPOs up to an energy of 30 keV and energy dependence of the RMS variability of the QPOs (Fig. 6) shows that the Keplerian frequency model is unlikely to be applicable in the case of XTE J1858 + 034. Considering the strong energy dependence of the QPO rms detected both during the 1998 and the 2004 outbursts, and the possibility of a correlation between the QPO frequency and the overall X-ray intensity as discussed above, the beat frequency model seems to be more suitable in the case of XTE J1858 + 034.

Spectral parameters of XTE J1858 + 034 obtained from the RXTE-PCA observations during the second outburst in 2004 are similar to those measured during the previous outburst (Paul & Rao 1998). The iron emission line has an equivalent width in the range of 125 eV to 200 eV and the line centre energy is in the range of 6.3–6.7 keV. With the moderate resolution spectral data from the RXTE-PCA

detectors we cannot conclude about the emission line characteristics, but it is likely that the emission line is from neutral or lowly ionized iron in the circumstellar material. The column density measured during the present observations is in the range of $(7.7\text{--}12.3) \times 10^{22}$ atoms cm^{-2} , which is higher than the same during the previous outburst. As the absorption column density depends on the mass loss history of the companion star prior to the outburst, a change in the column density is quite likely. We note that the iron line equivalent width and the absorption column density is consistent with a scenario in which the iron line is produced in a spherical shell of material which also causes the low energy X-ray absorption.

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Luminosity Dependent Study of the High Mass X-ray Binary Pulsar 4U 0114 + 65 with ASCA

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Abstract. Here we report the spectral characteristics of the high and low states of the pulsar 4U 0114 + 65 and examine the change in the parameters of the spectral model. A power law and a photoelectric absorption by material along the line of sight together with a high energy cut-off suffice to describe the continuum spectrum in both the states. A fluorescence iron line at ~ 6.4 keV is present in the high as well as in the low state, though it is less intense in the latter. The photon index, cut-off energy and e-folding energy values hardly show any discernible change over the states. We compare these spectral characteristics as observed with ASCA with those of other satellites. We also compare the spectral characteristics of 4U 0114 + 650 with other X-ray sources which show intensity variation at different time scales.

Key words. Stars: pulsar: individual (4U 0114 + 65)—X-rays: pulsar.

1. Introduction

4U 0114 + 650 is a pulsar which belongs to the class of High Mass X-ray Binaries (HMXB). It was discovered in the SAS-3 Galactic Survey (Dower & Kelley 1977). The pulsar has a B1Ia supergiant optical companion and the binary system is at a distance of ~ 7 kpc (Reig *et al.* 1996). Crampton *et al.* (1985) reported an orbital period of ~ 11.59 days for this HMXB from the optical radial velocity measurements. Their measurements were not able to distinguish between a circular or elliptical orbit for this binary. Corbet *et al.* (1999) obtained X-ray intensity modulations at a period of ~ 11.63 days from the long term light curve of RXTE-ASM. In addition to the orbital modulation, the X-ray light curve of this source shows considerable variability, with flaring activity for a few hours (Apparao *et al.* 1991; Finley *et al.* 1992) and also short-term flickering for minutes (Koenigsberger *et al.* 1983). The presence of a ~ 2.8 hr periodicity has been observed in the X-ray light curves of 4U 0114 + 65 from the analysis of archival EXOSAT and ROSAT data (Finley *et al.* 1992). This has been re-confirmed by Corbet *et al.* (1999) with the data from RXTE-ASM and Hall *et al.* (2000) with RXTE-PCA. This periodicity has been interpreted as the spin period of the pulsar. Very recently, Farrell *et al.* (2005) reported the detection of a superorbital period of ~ 30.7 days in this pulsar by analyzing the data of RXTE-ASM.

The X-ray spectrum emanating from the pulsar is nicely fitted with a generic model applicable in the case of HMXBs consisting of a power law with an absorption along

the line of sight, modified at high energies by an exponential cut-off. There is also the presence of a fluorescent line at ~ 6.4 keV due to neutral or lowly ionized iron (Yamauchi *et al.* 1990; Masetti *et al.* 2005). The typical values of the model parameters are a photon index ~ 1 , the cut-off energy at ~ 8 keV and the e-folding energy at ~ 20 keV. The column density values vary from $\sim 3 \times 10^{22}$ atoms cm^{-2} (Hall *et al.* 2000, with RXTE) to $\sim 15 \times 10^{22}$ atoms cm^{-2} (Masetti *et al.* 2005, with BeppoSAX). X-ray spectroscopy of this source during the different intensity states has been carried out in different energy bands with data from various satellites. Like, Koenigsberger *et al.* (1983) with combined data from Einstein Observatory, HEAO-1 and OSO-8, Yamauchi *et al.* (1990) with GINGA, Apparao *et al.* (1991) with EXOSAT, Hall *et al.* (2000) with RXTE, Bonning & Falanga (2005) with INTEGRAL and Masetti *et al.* (2005) with BeppoSAX.

A very preliminary overall X-ray spectral analysis of this source observed by ASCA was reported earlier (Ebisawa 1997). The preliminary analysis showed the presence of a fluorescent iron-line at ~ 6.4 keV and reported a variable hydrogen column density (N_H). But the broad-band spectral characteristics of 4U 0114 + 65 has now been established with BeppoSAX which confirms the presence of a high energy cut-off in the spectra in both the states. Since the cut-off energy falls within the ASCA energy band, we have reanalyzed the ASCA data and determined the column density and iron-line parameters accurately. Since ASCA has a better spectral resolution compared to other detectors which have observed this source and it also has a low energy coverage comparable to BeppoSAX, it is imperative to investigate the source in detail with ASCA. In this paper, we map the spectral change of 4U 0114 + 650 with ASCA at two different intensity levels. We also compare the spectral characteristics of this HMXB with the results obtained from other satellites at different orbital phases. For this purpose, we have described the observations in section 2, the data analysis and results in section 3 and thereafter discuss our results in the last section.

2. Observations

The present observation of the pulsar 4U 0114 + 65 was carried out by ASCA on 1997-02-10 (at an orbital phase of 0.19) for a useful exposure time of ~ 25 ks. The time span between the start and the end of the observation was ~ 55 ks, which is about 5.5% of the orbital period. ASCA has two gas-imaging spectrometers (GIS) and two solid-state imaging spectrometers (SIS). The energy resolutions are 500 eV and 130 eV (FWHM) @ 6 keV for the GIS and SIS detectors respectively. For more details about ASCA, the reader is referred to Tanaka *et al.* (1994). During this observation, the GIS was operated in the standard PH mode and the SIS was used in the Bright mode.

3. Data reduction, analysis and results

For this work, we have used the standard data reduction procedure of the ASCA guest observer facility. We have taken the archival screened data for our analyses. The source light curves and the spectra were obtained from circular regions of radii of $6'$ and $3'$ around the centroid of the source from the GIS and SIS respectively. The background

light curves and spectra were extracted from the source-free part of the field-of-view from similar circular regions. For better statistics, the light curves and spectra from the GIS2 and GIS3 detectors, as well as the data from the SIS0 and SIS1 detectors were combined together. A detailed description of ASCA data reduction and analysis can be found in <http://heasarc.gsfc.nasa.gov/docs/asca/abc/abc.html>

3.1 The light curve

We have shown the total GIS and SIS background subtracted light curves with a binsize of ~ 100 s in Fig. 1. The light curve shows flaring state during the first 20 ks and low state afterwards. On the average, there is a decay of the count rate from 10 ks onwards as can be seen from the light curves. The intensity at the peak of the flare is ~ 12 – 15 times than the persistent low level emission. At times, the intensity almost drops down to zero (at ~ 22 and ~ 32 ks in Fig. 1). In addition, there are short time-scale intensity variations for a few minutes. For our analyses, we have divided the whole light curve into two sections. The first section corresponds to the high state and is between 0 and 18 ks. The other half is ascribed to the low state of the source. The GIS and SIS average count rates (background subtracted) for the whole observation are 3.9 and 3.0 count s^{-1} respectively. The background subtracted count rates for the high state GIS and SIS light curves are 7.8 and 6.5 count s^{-1} respectively, whereas the low state count rates for the said instruments are 2.0 and 1.7 count s^{-1} . We were not able to detect the ~ 2.7 hr periodicity from the light curves. Strong aperiodic variability at short and long time scales and inadequate exposure prohibits us from finding pulsations at the reported 2.7 hr spin period.

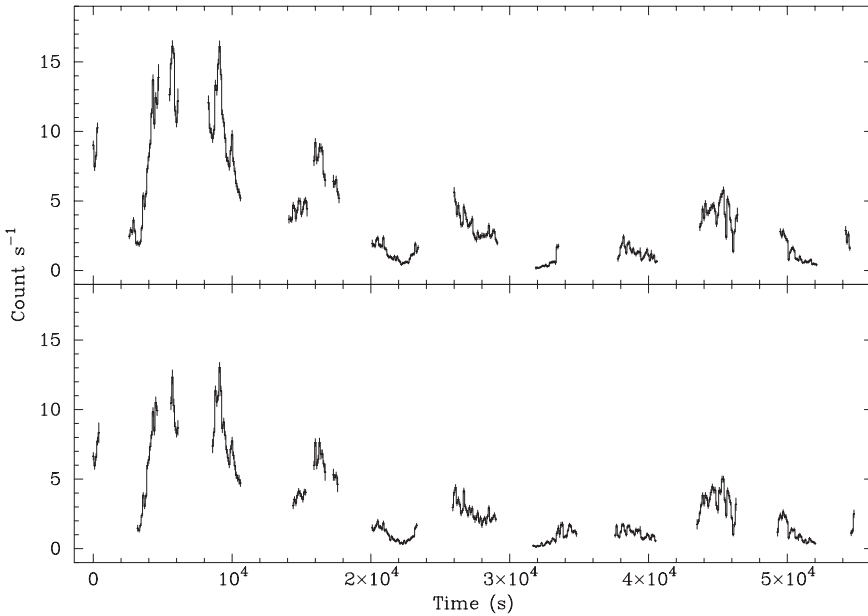


Figure 1. The background subtracted light curves of 4U 0114 + 65 taken on 1997-02-10 with detectors onboard the ASCA Observatory are shown here with binsize of 100 s. The top panel shows the GIS light curve in the 0.7–10.0 keV band and the bottom panel shows the SIS light curve in the 0.5–9.0 keV band.

3.2 The spectral model and the fitting

The spectral fitting was carried out in XSPEC version 11.2.0 (Shafer *et al.* 1989). After appropriate background subtraction, the GIS and SIS spectra were fitted simultaneously. The energy bands used for the fitting was 0.7–10.0 keV for the GIS and 0.5–9.0 keV for the SIS, where the effective areas and energy responses of the detectors are well defined. The 1024 channels of the GIS and the 512 channels of the SIS spectra were grouped suitably. In all of our fits described below, we have included an interstellar photoelectric absorption along the line of sight (*wabs* in XSPEC). A constant factor was incorporated to account for the relative normalization of the GIS and SIS detectors.

We extracted the spectra for the high and the low intensity levels (as defined above) respectively. The high state spectrum was first fitted with a simple power law and residuals were seen between ~ 6.2 and 6.6 keV. This prompted us to include a Gaussian line in the model. And that expectedly improved the fit from a χ^2 of 417 for 265 degrees of freedom to a χ^2 of 318 for 262 degrees of freedom. Though the reduced χ^2 of 1.2 was reasonably acceptable, there apparently remained some residuals between 1.0 and 3.0 keV. Also from ~ 7 keV onwards there was a hint of a high energy cut-off in the residuals. Hence we added the high energy cut-off in our model. That helped to weed out the above mentioned features in the residuals altogether. Figure 2 shows the high state spectrum along with the residuals. The low state spectra was fitted likewise and we obtained an acceptable fit without any feature in the residuals. The low state spectrum is shown in Fig. 3 with the best fit model and the residuals.

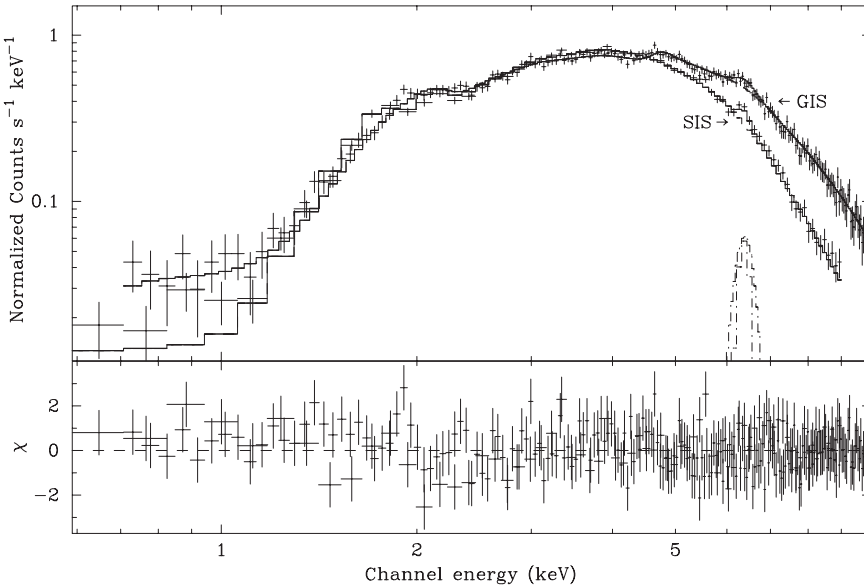


Figure 2. The high intensity level X-ray spectrum of 4U 0114 + 65 measured with the ASCA SIS and GIS are shown in the top panel along with the best fitted spectral model components as histograms. Contribution of each data point in the spectrum to the total χ^2 is shown in the bottom panel.

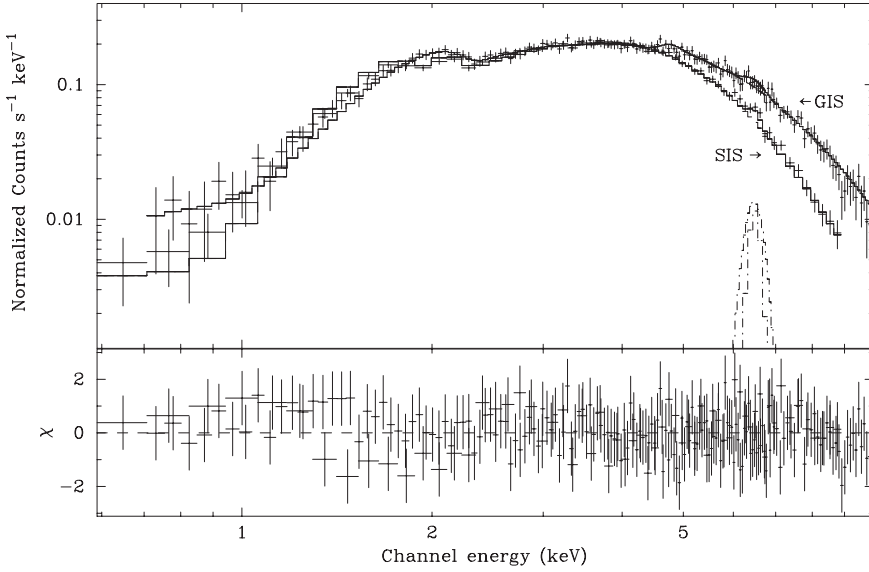


Figure 3. Same as in Fig. 2 but for the spectrum at low intensity level.

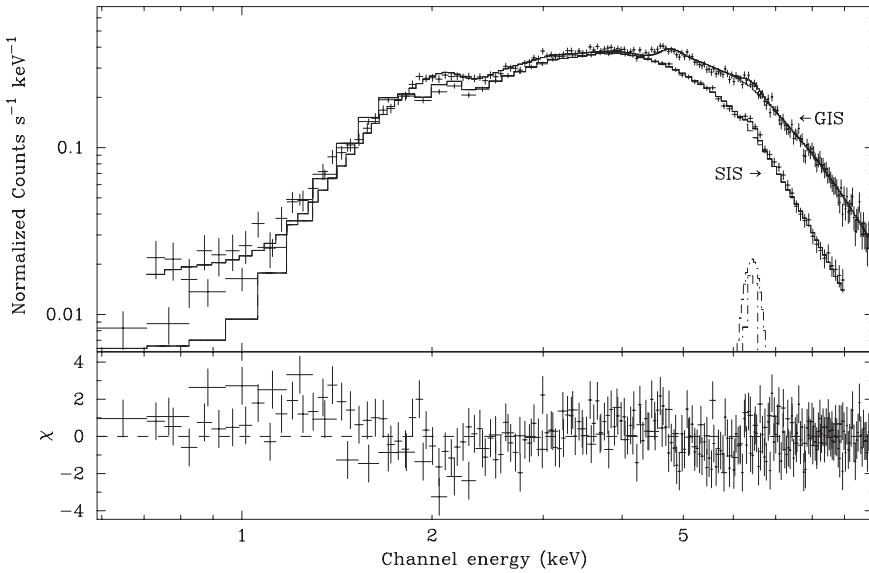


Figure 4. The spectrum measured during the entire observation is shown here with all other details same as in Fig. 2. Unlike the spectra in the low and high intensity levels (Figs. 2 and 3), the residuals show systematic differences between the data and the best fitted spectral model.

After fitting the high and low state spectra as above, we tried to fit the total spectrum (for the full time duration) with the same model. Though we obtained a reasonable reduced χ^2 of ~ 1.2 , the parameters, except the line-centre energy, could not be constrained and the best fit model shows structured residuals between 1.0 and 3.0 keV (Fig. 4). In Table 1, we give the best fit model parameters with the 90% error estimates.

Table 1. The spectral parameters of the states with 90% error bars (N_H : Column density, PL norm: Power law normalization and $L_{2-10 \text{ keV}}$: Unabsorbed luminosity in the 2–10 keV range).

Parameters	High	Low	Total
N_H (10^{22} cm^{-2})	$3.44^{+0.13}_{-0.12}$	$2.01^{+0.14}_{-0.18}$	2.81
Γ (Photon index)	$0.64^{+0.05}_{-0.05}$	$0.49^{+0.11}_{-0.15}$	0.62
PL norm (photons $\text{keV}^{-1} \text{ cm}^{-2} \text{ s}^{-1}$)	$0.03^{+0.002}_{-0.002}$	$0.005^{+0.001}_{-0.001}$	0.013
Line energy (keV)	$6.38^{+0.03}_{-0.06}$	$6.49^{+0.08}_{-0.13}$	$6.43^{+0.07}_{-0.06}$
Line width (keV)	$0.006^{+0.11}_{-0.006}$	$0.013^{+0.24}_{-0.003}$	0.016
Line flux ($10^{-4} \text{ photons cm}^{-2} \text{ s}^{-1}$)	$5.60^{+11.76}_{-1.55}$	$1.63^{+9.16}_{-0.70}$	2.07
Equivalent width (eV)	64.5	73.0	49.5
Cut-off energy (keV)	$6.47^{+0.38}_{-0.35}$	$4.41^{+0.37}_{-0.28}$	6.17
e-folding energy (keV)	$8.54^{+4.45}_{-1.65}$	$7.82^{+0.84}_{-1.04}$	7.83
Red. χ^2 (d.o.f.)	0.86 (260)	0.65 (219)	1.18 (260)
$L_{2-10 \text{ keV}}$ ($10^{36} \text{ ergs s}^{-1}$)	3.65	0.80	1.71

For the composite spectrum, the model parameters are not constrained well and error estimates are not given. Continuum parameters are identical within error-bars while the line flux is lower by a factor of ~ 4 . The e-folding energy is not well constrained in the high state. This is possibly due to the limited energy range offered by ASCA. But in the low state, it is comparatively better constrained.

4. Discussion

The main objective of our work was to measure the spectral parameters of the pulsar 4U 0114 + 65 in its high and low states with the now known model components that fit the broad band spectrum well and to study the iron emission line parameters *vis-à-vis* the column density (N_H) and continuum parameters at different orbital phases and intensity levels. As enunciated in the results, the most suitable model describing the source in the different intensity states with ASCA was found to be a power law modified by a cut-off, along with an exponential absorption along the line of sight. The presence of a Gaussian line at $\sim 6.4 \text{ keV}$, the fluorescent iron line, was also confirmed in both the spectra. It can be seen that as the unabsorbed luminosity between the two states varies by a factor of ~ 4.5 , there is a little change in the spectral parameters of the continuum, i.e., in the photon index and the column density (N_H). The most significant difference between the high and low state spectra is a change in the iron-line flux by a factor similar to the change in the continuum.

It is pertinent to compare our results with those obtained from other satellites. In order to do that, we have projected in a nutshell in Table 2, the spectral parameters of this pulsar as measured with other satellites. We have also mentioned the orbital phases during which those observations were made to examine if there is any orbital phase dependence of the spectral parameters. The orbital period and the epoch was taken from Corbet *et al.* (1999). We have specifically mentioned the high and low states in the table to bring out any possible differences.

Table 2. The spectral parameters obtained from various satellites. The high and low states being designated as H and L respectively. (N_H : Column density, Γ : Power law index, E_c : Cut-off energy, E_f : e-folding energy, EW: Equivalent width of the 6.4 keV Fe line, L: Luminosity) and (ϕ : Orbital phase, S: Satellite, A: ASCA, B: BeppoSAX, I: Integral, R: RXTE, EX: EXOSAT, G: GINGA, EN: Einstein).

ϕ (S)	N_H (10^{22} cm^{-2})	Γ	E_c (keV)	E_f (keV)	EW (eV)	L ($10^{36} \text{ ergs s}^{-1}$)
0.19 (A)	$3.44^{+0.13}_{-0.12}$	$0.64^{+0.05}_{-0.05}$	$6.47^{+0.38}_{-0.35}$	$8.54^{+4.45}_{-1.65}$	64.5	3.65 (H)
	$2.01^{+0.14}_{-0.18}$	$0.49^{+0.11}_{-0.15}$	$4.41^{+0.37}_{-0.28}$	$7.82^{+0.84}_{-1.04}$	73.0	0.80 (L)
0.67 (B)	$9.7^{+0.7}_{-0.9}$	$1.33^{+0.09}_{-0.16}$	$12^{+2.0}_{-3.0}$	$21^{+4.0}_{-3.0}$	—	1.35 (H)
	$15.4^{+2.0}_{-1.7}$	$0.9^{+0.2}_{-0.2}$	$6.0^{+0.9}_{-0.7}$	$17^{+5.0}_{-3.0}$	80	0.37 (L)
0.85 (I)	—	$1.6^{+0.5}_{-0.5}$	$9.0^{+14.1}_{-8.8}$	$22.1^{+12.1}_{-6.0}$	—	1.8 (H)
0.82 (R)	$3.2^{+0.3}_{-0.3}$	$1.28^{+0.04}_{-0.05}$	$8.1^{+0.4}_{-0.4}$	$19.2^{+1.1}_{-1.0}$	—	1.25 (H)
	$4.1^{+0.4}_{-0.4}$	$1.63^{+0.04}_{-0.02}$	$11.3^{+0.8}_{-0.9}$	$22.4^{+6.7}_{-4.7}$	250	0.2 (L)
0.35 (EX)	$9.4^{+1.9}_{-1.9}$	$2.47^{+0.30}_{-0.30}$	—	—	526	0.45 (L)
0.36 (G)	$2.9^{+0.2}_{-0.2}$	$0.92^{+0.02}_{-0.02}$	$7.1^{+0.3}_{-0.3}$	$16.1^{+0.8}_{-0.8}$	70	5.0 (H)
	$4.7^{+0.6}_{-0.6}$	$1.07^{+0.03}_{-0.03}$	$6.8^{+1.0}_{-1.0}$	$15.4^{+4.0}_{-4.0}$	340	0.33 (L)
0.78 (EN)	$3.2^{+1.0}_{-1.0}$	$1.2^{+0.1}_{-0.1}$	—	—	—	2.11 (H)
	$3.2^{+1.0}_{-1.0}$	$1.2^{+0.3}_{-0.3}$	—	—	—	0.12 (L)

Note: The luminosities quoted in both the tables have been calculated with the distance to the pulsar as 7 kpc.

The RXTE and GINGA observations found marginal evidence for an increase of N_H during the low states while the BeppoSAX observations found N_H to increase by a factor of 2 in the low state. We note that in any of the three cases mentioned above, the increase in N_H by itself cannot explain occurrence of the low state. The galactic column density along the line of sight of 4U 0114 + 65 is $\sim 0.8 \times 10^{22} \text{ cm}^{-2}$. This is about an order of magnitude lower than the column density values of 4U 0114 + 65 obtained previously or in the present work. The N_H values measured with ASCA seem to compare well with that obtained with RXTE and GINGA, though the ASCA observation shows a slight decrease in the value of N_H in the low state, and the N_H values in the present work are a factor of ~ 2 –4 times lesser than that measured with BeppoSAX. From Table 2, if we compare only the high state observations or only the low state observations, we do not find any clear orbital phase dependence of the N_H or the equivalent width of the iron line. If the soft X-ray absorption is due to the material in the outflowing wind from the companion star and if the iron emission line is produced by the reprocessing of the continuum hard X-rays in this material, a smooth orbital phase dependence of the N_H and the equivalent width on orbital phase is expected. A similar finding in another pulsar GX 301–2 was interpreted in terms of a clumpy structure of the stellar wind (Mukherjee & Paul 2004).

Regarding photon index of the power law component, Table 2 shows that apart from the results from EXOSAT which point towards a very steep power law, the results from other satellites generally indicate a Γ of ~ 1 . The unusually high value of Γ with

EXOSAT may be due to the slightly different spectral model used and the limited band width. The values of Γ obtained with ASCA is somewhat lower than 1. On the whole it is seen that Γ for this pulsar does not change appreciably with X-ray intensity.

From Table 2, we notice that the equivalent width of the fluorescent iron-line is comparatively higher in some of the low state observations and on most occasions (with BeppoSAX and RXTE), the 6.4 keV line is better detected in the low state. However with ASCA, we report the detection of the iron-line in both the states with appreciable and similar equivalent widths (Table 2). Since ASCA has the best spectral resolution in comparison to the other detectors mentioned herein, the measurements of the iron line parameters is likely to be most accurate with ASCA.

A hint of a soft excess was reported in the BeppoSAX spectra of this source by Masetti *et al.* (2005). They showed an apparent presence of the soft excess below 3 keV which they fitted with a ~ 0.3 keV thermal component. However, in any of the ASCA fits covering an energy band lower compared to the reported BeppoSAX spectrum, we did not find any signature of the so-called soft excess as found in the spectra of many X-ray pulsars (Paul *et al.* 2002; Hickox *et al.* 2004). Moreover, in the case of 4U 0114 + 65, for the unabsorbed flux observed with ASCA in the high state ($\sim 6 \times 10^{-10}$ ergs cm $^{-2}$ s $^{-1}$) along with the N_H of 3.4×10^{22} atoms cm $^{-2}$ falls into the region in Fig. 1 of Hickox *et al.* (2004) which shows a collection of pulsars lacking soft excess.

During our observation, the transition of 4U 0114 + 650 between the high and the low state is abrupt. Similar abrupt flux changes have been observed in another pulsar GX 1 + 4 which also shows aperiodic transitions between its high and low states and thus requires special mention here. Naik *et al.* (2005) studied GX 1 + 4 for an extended period with observations by BeppoSAX and reported an order of magnitude decrease in the X-ray flux for a part of the observation, similar to the present observation of 4U 0114 + 650. In GX 1 + 4, the absorption column density was an order of magnitude higher in the low state, which is not the case for 4U 0114 + 650. In addition, Naik *et al.* (2005) found a strong increase in the iron line equivalent width in the low state which is also different from what we have observed in 4U 0114 + 650.

Equivalent width of the iron line is found to increase appreciably in the low states of several other pulsars like Her X-1 (Naik & Paul 2003), LMC X-4 (Naik & Paul 2004) and SMC X-1 (Vrtilek *et al.* 2005). The high and low states of the latter sources are part of the super-orbital period arising due to a precessing warped inner accretion disk. Though the pulsar 4U 0114 + 650 is reported to have a super-orbital period (Farrell *et al.* 2005), the abrupt decrease in X-ray flux over a few thousand second time scale reported here can not be assigned to the precession of a warped accretion disk. The change in luminosity during our observation can be ascribed to a local change in the density of the stellar wind from which the pulsar accretes material. The rise and fall times of 1–2 ks as can be seen in Fig. 1, and an assumed orbital velocity of a few hundred km per sec indicates clumpiness of the stellar wind at a length scale of about 10^{10-11} cm. The long term light curve of this source measured with the RXTE-ASM does not have enough sensitivity to measure incoherent intensity variations at the time scale of a few ks. During the ASCA observation, the iron line flux decreased with luminosity and has a direct correlation with the continuum luminosity (Table 1). It supports a picture in which the amount of fluorescence material, like in the form of a circumstellar shell at a large distance has remained unchanged but the local accretion has changed.

5. Conclusion

In this paper, we have reported an accurate measurement of the spectral parameters of the HMXB pulsars 4U 0114 + 650 in two different intensity levels. We have detected iron emission lines of similar equivalent width in both the intensity levels and the other spectral parameters are found to be similar in the two intensity levels. This shows that the intensity variations at a time scale of a few hours observed in 4U 0114 + 650 was not due to increased absorption, but possibly due to wind density that causes changes in the accretion rate. We have compared the present observation of this object with that done by other X-ray observatories and also with other HMXB pulsars which show widely different intensity states. 4U 0114 + 650 is found to have characteristics different from many other HMXB pulsars.

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Photometric Modelling of Close Binary Star CN And

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Abstract. The results of two color photometry of active close binary CN And are presented and analyzed. The light curves of the system are obviously asymmetric, with the primary maximum brighter than the secondary maximum, which is known as the O'Connell effect. The most plausible explanation of the asymmetry is expected to be due to spot activity of the primary component. For the determination of physical and geometrical parameters, the most new version of W-D code was used, but the presence of asymmetry prevented the convergence of the method when the whole light curves were used. The solutions were obtained by applying mode 3 of W-D code to the first half of the light curves, assuming synchronous rotation and zero eccentricity. Absolute parameters of the system were obtained from combining the photometric solution with spectroscopic data obtained from radial velocity curve analysis. The results indicate the poor thermal contact of the components and transit primary minimum. Finally the O-C diagram was analyzed. It was found that the orbital period of the system is changing with a rate of $dP/dt = -2.2(6) \times 10^{-10}$ which corresponds to mass transfer from more massive component to less massive with the rate of $dM/dt \sim 4.82 \times 10^{-8} M_{\text{sun}}/\text{year}$.

Key words. Close binaries—stars spot—active stars—light curve analysis.

1. Introduction

The variability of CN And (= BD + 39 59) was discovered by Hoffmeister (1949). On the basis of photographic data, Tsesevich (1956) classified the system as an Algol-type binary and derived an orbital period of 2.599 days. Later Lochel (1960) classified it as a WUMa-type system and was the first to correctly determine the period of 0.462798 days. The system has been observed by many authors (e.g., Seeds and Abernathy 1982; Michaels *et al.* 1984; Evren *et al.* 1987). The light curves are characterized by interesting asymmetries (Rafert *et al.* 1985) similar to V1010 Ophiuchi binaries (Shaw 1994). Its light level of activity was attested by two flares seen by Yu-Lan and Quing-Yao (1985), as well as its X-ray luminosity of $\log Lx = 30.55$ (Shaw *et al.* 1996). Two groups of investigators have attempted analyzing the light curves for geometrical and

Table 1. Observed times of minima of CN And.

J.D.Hel 2451000+	Type of Min	(O-C) ₁	(O-C) ₂
458.3786	I	-0.0250	-0.0058
469.2556	II	-0.0236	-0.0043
471.3341	I	-0.0277	-0.0083
807.3239	I	-0.0262	-0.0045
811.2607	II	-0.0231	-0.0014
814.2692	I	-0.0227	-0.0010

physical parameters, disregarding the lack of any information on spectroscopic mass ratio. Kaluzny (1983) expected the mass ratio to be within $0.55 < q_{\text{ph}} < 0.85$. Rafert *et al.* (1985) found the most likely interval to be $0.5 < q_{\text{ph}} < 0.8$; they saw indications of strong contact. In the present work, the light curves of the system CN And have been analyzed on the basis of new photoelectric observations carried out by authors and spectroscopic mass ratio given by Rucinski *et al.* (2000).

2. Observations and data

The present observations of CN And were carried out with 40 cm Cassegrain telescope of Khadjeh Nassir Addin Observatory of Tabriz University during the summer of 1999 and 2000. The observations were made in B and V filters with peak wavelength (widths) of $4350 \text{ \AA}$ ($970 \text{ \AA}$) and $5500 \text{ \AA}$ ($850 \text{ \AA}$) respectively. A single channel photometer equipped with an unrefrigerated photomultiplier tube (RCA 1P21) was used. The output current was amplified and fed into a PC via an A/D converter for rapid data access. Two stars GSC 02787-01780 and BD + 39 65 were used as comparison and check stars respectively. During the observations four primary and two secondary times of minima were obtained in each filter, by parabolic fitting to the observational points, making use of the least square method. Those minima are given in Table 1 where $(O-C)_1$ and $(O-C)_2$ residuals have been computed using equations

$$\text{J.D.Hel (Min I)} = 2450698.9591(18) + 0.46279372(13) \times E$$

and

$$\begin{aligned} \text{J.D.Hel (Min I)} = & 2450698.9447(14) + 0.46279092(19) \\ & \times E - 9.8(6) \times 10^{-11} \times E^2 \end{aligned}$$

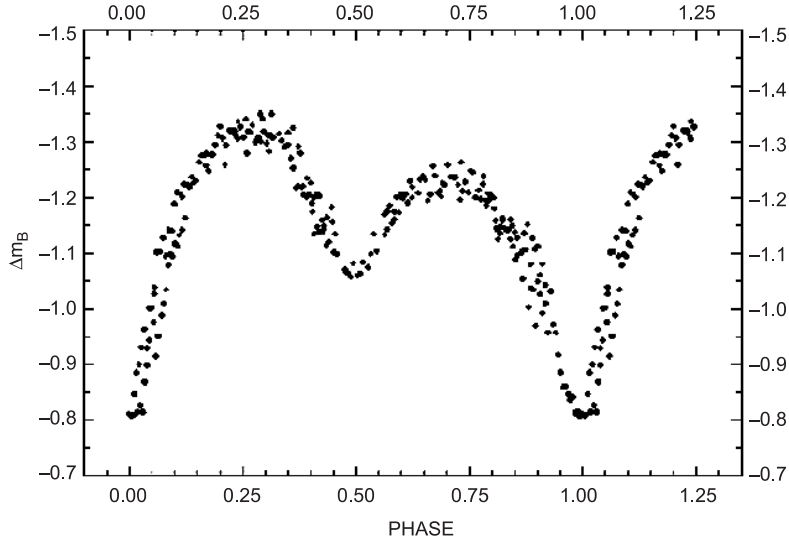
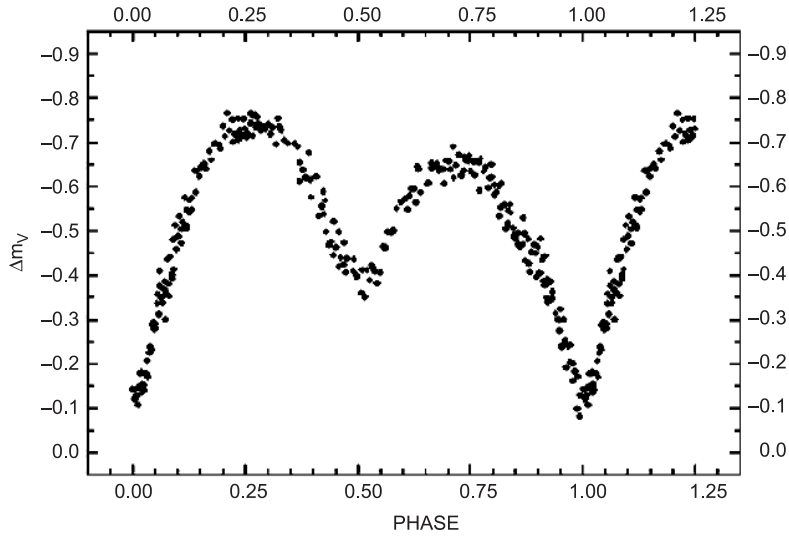
given by Samec *et al.* (1998) respectively.

We obtained ~ 300 individual points for each filter with standard error of single observation ~ 0.02 mag. The data were reduced to outside atmosphere and to standard system based upon comparison star reading and standard stars observations. Standard magnitudes and colors of variable stars in different phases are given in Table 2.

Figures 1 and 2 show the corrected light curves. The light curves are obviously asymmetric, with the primary maximum (phase 0.25) being brighter than the secondary maximum (phase 0.75) which is known as the O'Connell effect. $\Delta(\text{Max I} - \text{Max II})$ is ~ 0.09 in B and ~ 0.10 in V filters.

Table 2. Standard magnitudes and colors of CN And in different phases.

Phase	V (mag)	$(B-V)$
0.00	10.43 ± 0.02	0.47 ± 0.02
0.25	9.78 ± 0.01	0.43 ± 0.01
0.50	10.13 ± 0.02	0.42 ± 0.01
0.75	9.88 ± 0.01	0.47 ± 0.02

**Figure 1.** Standard ΔB (Var – Comp) light curve of CN And for observations in 2000.**Figure 2.** Standard ΔV (Var – Comp) light curve of CN And for observations in 2000.

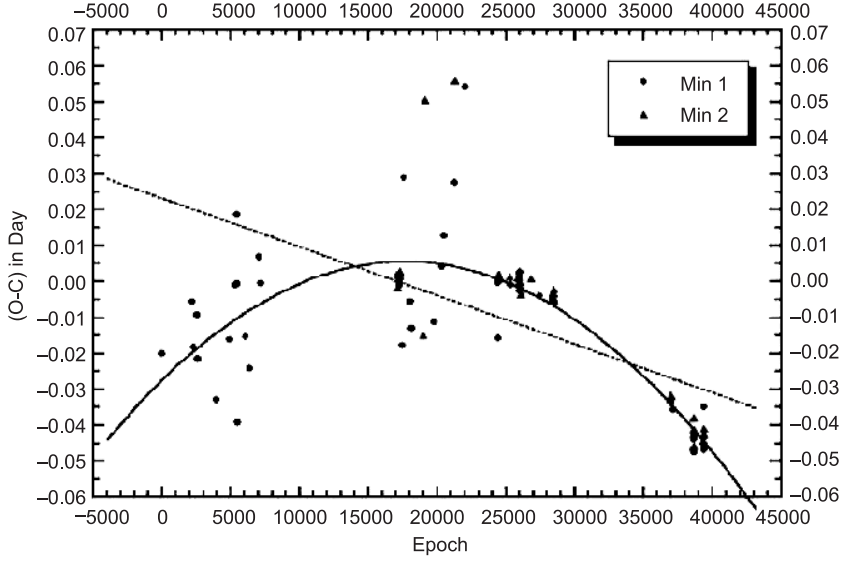


Figure 3. Linear and quadratic fit of all minima.

3. Period variation

From a comprehensive discussion of all times of minima of CN And, Samec *et al.* (1998) concluded that a representation of the observed times of minima could be made by a parabola, implying a steady rate of period variation of $dP/dt = -4.24 \pm 0.26 \times 10^{-11}$. We have re-investigated these data including also our own values given in Table 1.

We assigned weights of 1, 2, and 6 to visual, photographic and photoelectric minima respectively. Linear and quadratic weighted least-square fits to all minima (Fig. 3) yielded ephemerides:

$$\text{J.D.Hel (Min I)} = 2433570.5067(\pm 0.0026) + 0.462793293(\pm 0.000000076)\text{E}$$

and

$$\begin{aligned} \text{J.D.Hel (Min I)} = & 2433570.4567(\pm 0.0018) + 0.462798329(\pm 0.000000148) \\ & \times \text{E} - 1.048(\pm 0.0298)10^{-10} \text{E}^2. \end{aligned}$$

Residuals for each of these fits are shown in Figs. 4 and 5. The corresponding dP/dt is $-2.2(6)10^{-10}$. The significance of this period changes will be discussed in section 5.

4. Analysis

We used Wilson and Devinney (1971) differential correction method to analyze the observed light curves. It was assumed that CN And has a circular orbit and the rotational and orbital spins are synchronous. Also black-body models were employed. In deriving the photometric solution, the following parameters were adopted: Standard values of

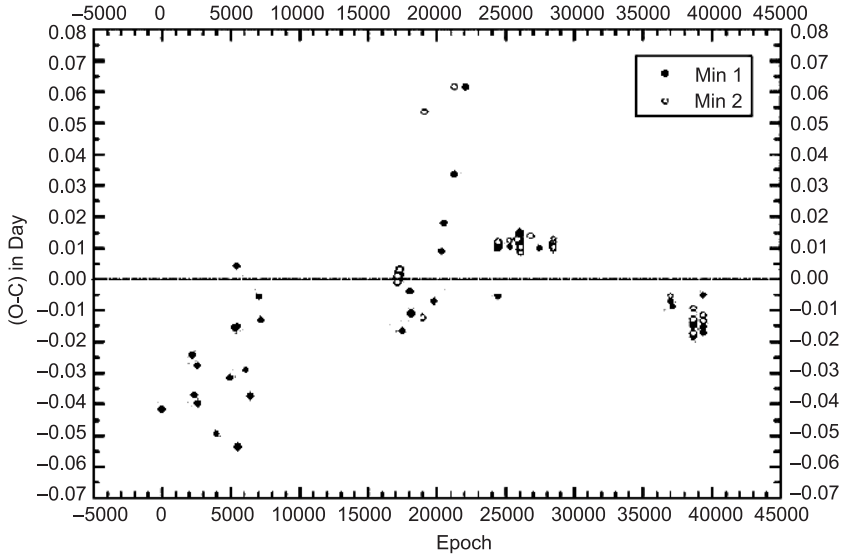


Figure 4. Residuals for linear least-square fit.

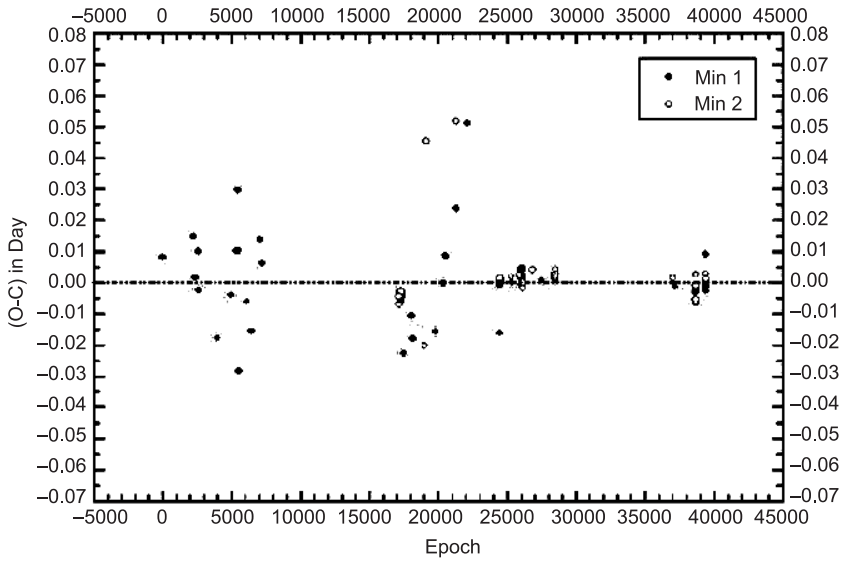


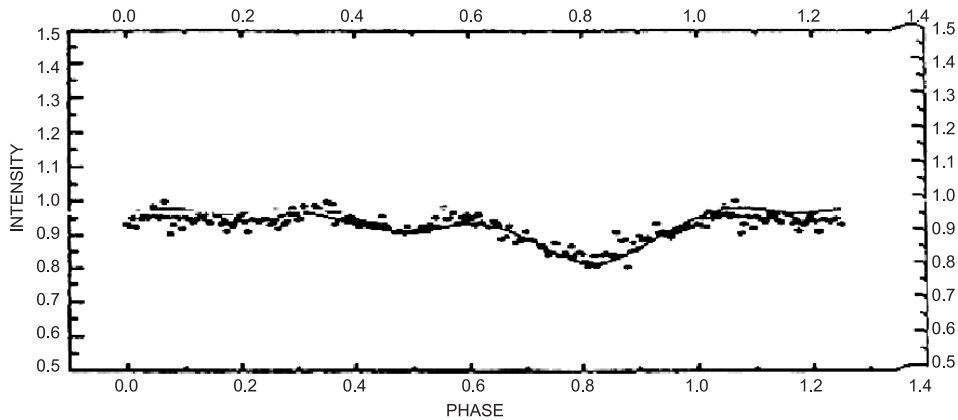
Figure 5. Residuals for quadratic least-square fit.

gravity darkening coefficients $g_1 = g_2 = 0.32$ for B and V filters and the bolometric albedoes $A_1 = A_2 = 0.5$ consistent with that of convective atmospheres (Lucy 1967). The limb darkening coefficients appropriate to spectral types of both components (F5 + G5) were found by interpolation from tables given by Al-Naimiy (1978).

Unfortunately, the information from light curves are not usually sufficient to derive an accurate value for mass ratio, q . Spectroscopic observations, on the contrary, lead to a more reliable and relatively accurate value for q . For the case of CN And, a mass ratio

Table 3. Spots - parameters.

Parameter	Spot no. 1	Spot no. 2
Co-latitude	70°	90°
Longitude	294°	174°
Angular-size	25°	18.5°
Temperature	5236° k	5544° k

**Figure 6.** Intensity variation due to spots and the theoretical curve fit based on parameters given in Table 3.

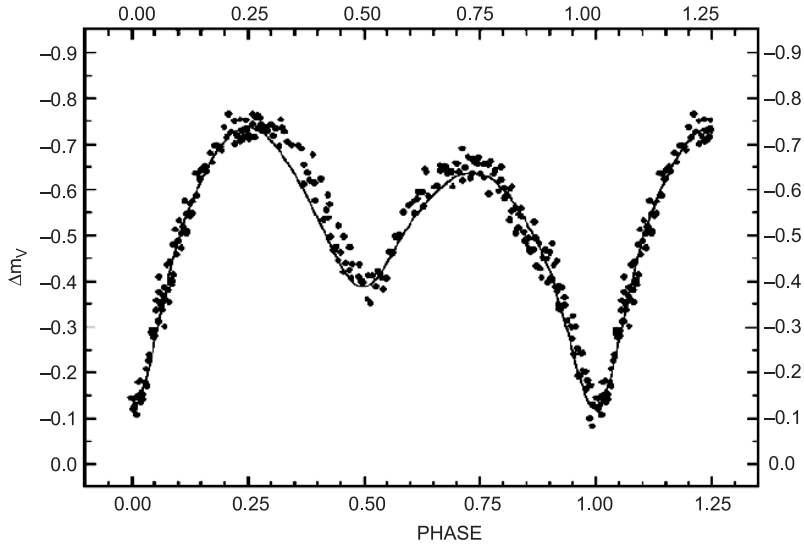
of 0.39 has been reported from radial velocity observations by Rucinski *et al.* (2000). We adopted this value for q . W-D code in mode 3 (contact and near contact configurations) was employed in which the luminosity of secondary component is coupled and cannot vary independently. The presence of asymmetry prevented the convergence of the method when the whole light curves were used.

The solutions were obtained in three steps as described below:

- Mode 3 of W-D code was applied to the first half of the light curves to obtain uncleaned parameters specifying the system.
- We deconvolved the theoretical light-curve corresponding to our adopted model in the first step from the observed light curve. The object of this was to obtain the intrinsic light variation of the system. The quasi-sinusoidal light variations were interpreted in terms of cool star-spots covering a significant fraction of stellar photosphere. Two circular spots were assumed on the equator of the primary, each with different values of ΔT (temperature difference between spots and surrounding photosphere). The corresponding sizes of area covered by spots were obtained using a method based on Budding's theory (1977). This procedure was repeated for B and V bands. The sizes of spotted area were plotted versus ΔT for each filter. The intersection point of these curves corresponded to spots ~ 1000 degrees cooler than the surrounding photosphere. Spots parameters are given in Table 3. Figure 6 shows the theoretical fit to the observed light variations due to spots.
- We cleaned the original observations by correcting for the presence of the distur-

Table 4. Photometric parameters of CN And.

Parameter	<i>B</i>	<i>V</i>	Mean values
	$\lambda = 4350 \text{ \AA}$	$\lambda = 5550 \text{ \AA}$	
i	70.52 ± 0.17	69.22 ± 0.19	69.87
q	0.39	0.39	0.39
T_1	6138 ± 22	6182 ± 18	6160
T_2	4690 ± 11	4628 ± 13	4659
$\Omega_1 = \Omega_2$	2.625 ± 0.005	2.609 ± 0.007	2.617
$L_1/(L_1 + L_2)$	0.9263 ± 0.0013	0.9031 ± 0.0018	–
$L_2/(L_1 + L_2)$	0.0737	0.0969	–
r_1 (pole)	0.4322	0.4522	0.4422
r_1 (side)	0.4615	0.4869	0.4742
r_1 (back)	0.5017	0.5057	0.5037
r_2 (pole)	0.2816	0.2954	0.2885
r_2 (side)	0.3112	0.2918	0.3015
r_2 (back)	0.3415	0.3391	0.3403
Ω_{in}	2.657	2.657	2.657
Ω_{out}	2.419	2.419	2.419

**Figure 7.** Observed and theoretical light curve of CN And based on parameters given in Table 4 (V-filter).

tion wave, adding the theoretical distortion wave effects to the data with opposite signs. Mode 3 of W-D code was then used to find cleaned parameters to provide us with information about the physical properties of the system. The set of parameters for which a fairly good fit to the observed points was obtained, are summarized in Table 4. The theoretical light curves fit of CN And based on the parameters in Table 4 are shown in Figs. 7 and 8.

The geometric and physical parameters determined from photometric solution were

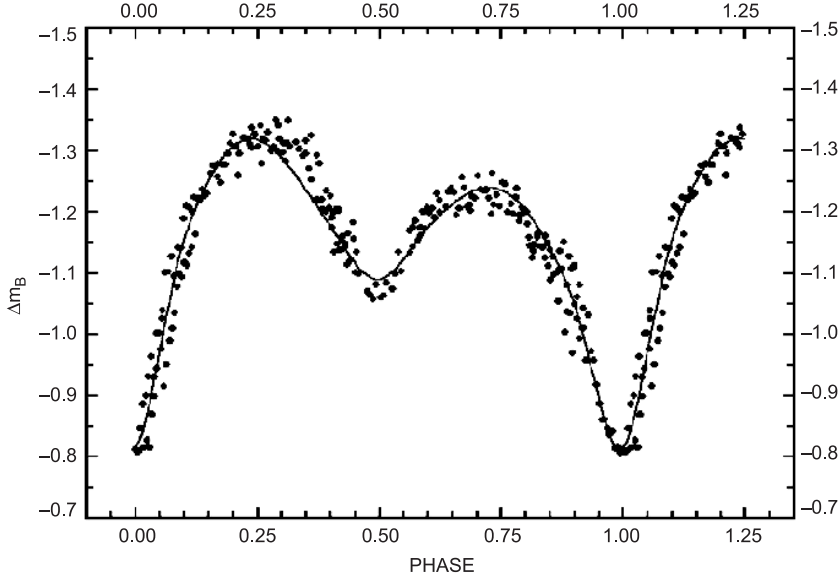


Figure 8. Observed and theoretical light curve of CN And based on parameters given in Table 4 (B-filter).

Table 5. Absolute dimensions of CN And calculated for mean values of Table 4.

Parameter	M/M_{sun}	R/R_{sun}	T/T_{sun}	L/L_{sun}
Primary	1.267	1.439	1.066	2.679
Secondary	0.494	0.943	0.806	0.376

combined with the spectroscopic data obtained from radial velocity analysis for both components by Rucinski *et al.* (2000) to yield the absolute dimension of the system (Table 5).

5. Concluding remarks

From the present study of the active binary CN And based on new photometric and spectroscopic data, we may conclude that:

1. The system CN And is an A-sub type WUMa binary with an over-contact configuration with fill-out factor defined as $f = (\Omega_{\text{in}} - \Omega_{1,2})/(\Omega_{\text{in}} - \Omega_{\text{out}})$ equal to 17%. This result is in agreement with that of Kaluzny (1983) and Rafert *et al* (1985) who found the system to be over-contact, but in contradiction to Van Hamme *et al.* (2001) who found semi-detached configuration.
2. The difference between the temperatures of the components is large (~ 1500 K), so the system is an over-contact binary with components in poor thermal contact.

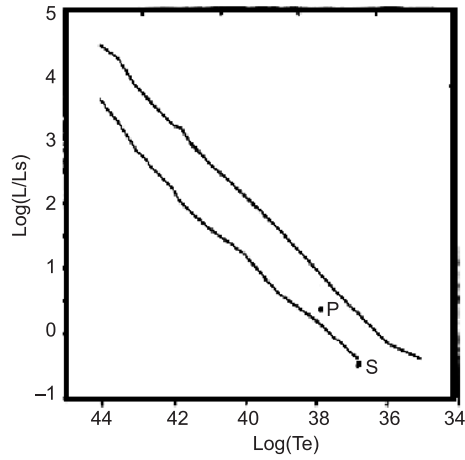


Figure 9. Positions of both components of CN And on H–R diagram.

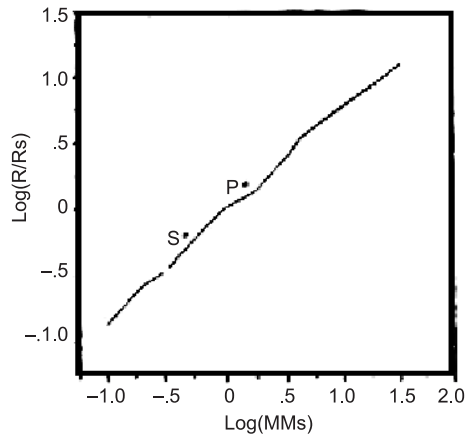


Figure 10. Positions of both components of CN And on R–M diagram.

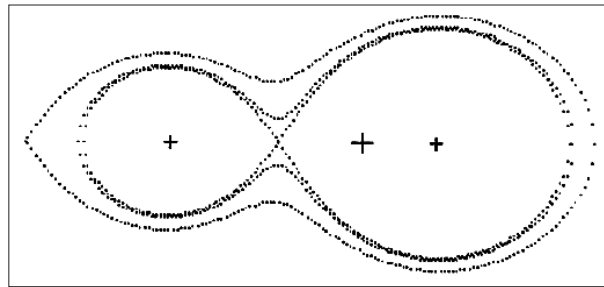


Figure 11. Configuration of the components of CN And in the orbital plane.

3. The positions of both components are shown in H–R and mass–radius diagrams in Figs. 9 and 10 respectively. It can be seen that the radii of both components are larger than the corresponding values for ZAMS stars with the same masses. So CN

And may be classified as an evolved over-contact undergoing thermal relaxation oscillation. Configuration of the system is shown in Fig. 11.

4. The period changes of the system (see section 3) which is decreasing continuously, could be interpreted in terms of mass transfer between two components. Assuming mass conservation, the rate of mass transfer is:

$$\Delta m/m = 3.472 \times 10^{-11}$$

or

$$dM/dt = 4.82(6) \times 10^{-8} M_{\text{sun}}/\text{year}.$$

Acknowledgements

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Preface

Transient phenomena occur on the Sun on many time scales putting out flashes of electromagnetic radiation and material ejection superposed on the thermal radiation and solar wind. The discovery of coronal mass ejections (CMEs) in the early 1970s and their connection to the interplanetary transient phenomena observed by space-borne instruments have dramatically improved our understanding of the Sun–Earth connected system. CMEs represent the most energetic aspects of solar eruptions that affect not only Earth’s space environment but also other planets in the solar system and ultimately the boundary of the heliosphere. The famous October–November 2003 events known as the Halloween CMEs amply demonstrated this by producing measurable effects throughout the heliosphere. CMEs are also considered to be the main contributor to the solar energetic population observed in the interplanetary medium by driving fast mode shocks that accelerate the particles. Direct impact of the CME plasma with its enhanced magnetic field is responsible for the large geomagnetic storms that have a number of consequences in geospace and on Earth’s surface. Although we understand the solar-terrestrial connection via the mass and photonic emissions from the Sun, we are far from when the eruptions occur on the Sun and how they evolve while propagating through the interplanetary medium. For example, we understand the CMEs as a magnetic phenomenon on the Sun but we do not have direct measurement of magnetic fields in CMEs. We have to work mainly with photospheric magnetic field measurements. The small-scale transients such as nanoflares, on the one hand, are related to the steady emissions from the Sun via coronal heating and solar wind acceleration. High speed solar wind from coronal holes interacting with neighboring slow wind form the corotating interaction regions (CIRs), which constitute another source of geomagnetic storms. Although the CIR-related storms are of smaller magnitude, they are more frequent and related to other effects such as production of relativistic electrons in the magnetosphere.

The set of papers in this special issue grew out of the presentations at the International Solar Workshop on Transient Phenomena on the Sun held at the Aryabhata Research Institute of Observational Sciences (ARIES) during April 5–7, 2005. This workshop was a part of the Golden Jubilee celebrations of ARIES, recognizing its outstanding contributions to solar and solar terrestrial physics. The workshop brought together Indian and international solar-terrestrial researchers for a three-day deliberation on issues related to solar and solar-terrestrial physics. The importance of international collaboration in data acquisition, analyses, theory and modeling is well recognized as evident from the growth of International Space Weather and Living with a Star communities in recent years.

This special issue has a large concentration of papers on the origin of CMEs and flares, and their space weather consequences. The open data policy of the solar and heliospheric observatory (SOHO) mission has fueled many of the reported studies. There are also papers on the quiet solar atmosphere and solar wind and their connection

to the solar interior and the dynamo in particular. India has a long tradition of solar-terrestrial studies over the last century. Continuation of this tradition is reflected in the instrumentation papers including those discussing the extension of the observational capability to space. The papers also reflect the new openness in data policy of the international community that has promoted enhanced scientific return from various data bases constructed from ground- and space-based observations.

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Why Study the Sun?

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Abstract. In this presentation we briefly describe the Sun through large number of illustrations and pictures of the Sun taken from early times to the present day space missions. The importance of the study of the Sun is emphasized as it is the nearest star which presents unparalleled views of surface details and numerous phenomena. Our Sun offers a unique celestial laboratory where a large variety of phenomena take place, ranging in temporal domain from a few milliseconds to several decades, in spatial domain from a few hundred kilometers to thousands of kilometers, and in the temperature domain from a few thousand degrees to several million degrees. Its mass motion ranges from thousandths to thousands of kilometers per second. Such an object provides us with a unique laboratory to study the state of matter in the Universe.

The existing solar ground-based and space missions have already revealed several mysteries of the outer environment of our Sun and much more is going to come in the near future from planned new sophisticated ground-based solar telescopes and Space missions.

The new technique of helioseismology has unravelled many secrets of the solar interior and has put the Standard Solar Model (SSM) on firm footing. The long-standing problem of solar neutrinos has been recently sorted out, and even the ‘back side’ view of the Sun can be seen using the technique of holographic helioseismology.

Key words. Sun—flares—prominences—coronal mass ejection—coronal holes—helioseismology.

1. Introduction

Let us bring our 4th magnitude, insignificant little star near enough to blaze into a minus 27th magnitude ball of fire – our Sun. This star is so near to us that we can see its minutest surface details. Because its warmth keeps this planet alive, it is the most important star for mankind. Every detail is worthy of observations and study.

In all ancient cultures, the Sun has occupied a central position. In many cultures and religions our Sun has been and is still being worshipped as the most powerful god and goddess. Large numbers of mythological stories have been woven around the Sun. The importance of the Sun had been well realized by the ancient man. It was used as a daily, seasonal, and annual clock. Huge monolithic structures like Stonehenge in England, Newgrange tunnel tomb in Ireland, Sun temples in Egypt, India and in many

other countries were built to observe, time, and predict important solar events like the summer and winter solstices, equinoxes, etc.

With the beginning of the telescope era, the Sun was one object in the sky that attracted the attention of many great astronomers like Galileo, Herschel, etc. Even before the telescope era, the Sun was observed by the naked eye for many centuries by Chinese and early Greek astronomers. More than 125 sightings of dark spots – *blemishes* on the Sun have been reported. There is even a Chinese reference of observing a ‘Black-blue vapour-cloud near an egg shaped dark spot’ (Clark & Stephenson 1978). I think the observer had indeed seen a white light flare and perhaps this is the first reported flare, even before the famous Carrington flare of 1859.

With the advent of telescopes, the solar observations and studies progressed by leaps and bounds. The discovery of solar rotation, the 11-year sunspot cycle, the tilt of the Sun’s axis by 7.5 degrees and many other discoveries followed. Very keen and persistent early solar observers like Father Secchi, Langley, Prof. Fernley and others, made extremely valuable solar observations. These observations showed remarkable details of solar features. For example in the left panel of Fig. 1 is shown a drawing of solar granulation and a sunspot pore made by Father Secchi in 1878, with an ordinary 13-inch refractor. Compare it with the modern picture taken with a state-of-the-art vacuum telescope equipped with image restoring systems as shown in the right panel of Fig. 1.

The right panel in Fig. 2 shows a drawing of a sunspot made by Father Secchi in 1878, displaying fine penumbral filaments and even the tiny sub arc second size umbral dots. How could he see such fine features and record them? In the left panel is shown a picture taken of a sunspot with the modern Dunn telescope at the NSO/Sac Peak Observatory, using adaptive optics.

With the introduction of physics and spectroscopy for the study of the Sun in the mid-nineteenth and early twentieth century, solar astronomy took another leap forward into the realm of solar astrophysics. Powerful new equipment, telescopes and spectrographs were located at suitable sites like Mount Wilson, Kodaikanal, Pic du

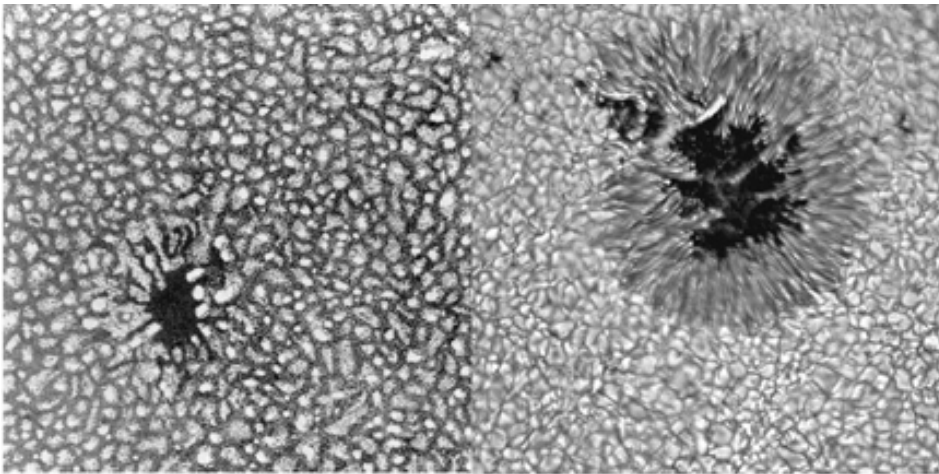


Figure 1. Left panel shows a drawing of solar granulations and a pore made visually by Father Secchi in 1878 with a 13-inch refracting telescope. Compare this with the high-resolution picture taken with modern day telescope shown in the right panel.

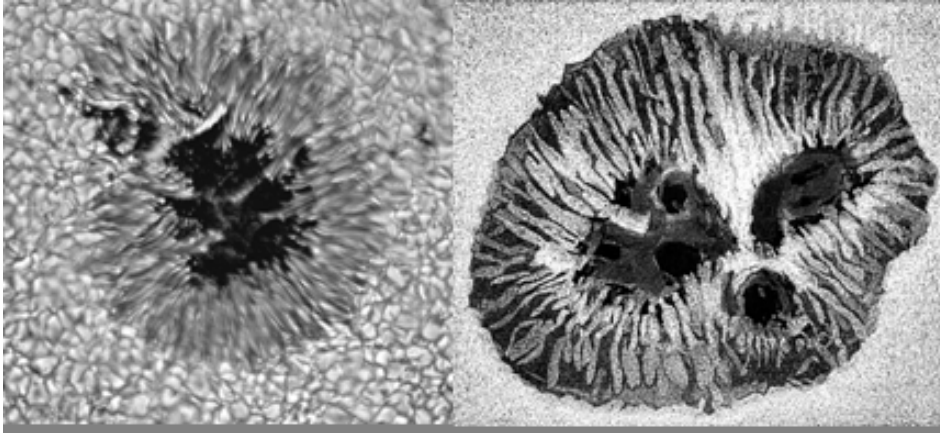


Figure 2. The left panel shows a picture taken of a sunspot with the modern Dunn telescope at the NSO/Sac Peak Observatory, using adaptive optics. In the right panel is shown a drawing of a sunspot made by Father Secchi in 1978, displaying fine penumbral filaments and even the tiny sub arc second size umbral dots.

Midi, etc., and making use of the new discoveries emerging in physics, helped to unravel several secrets of our Sun. Using spectroscopy, one important property of our Sun was revealed – that it is like an onion; you peel one layer after another and thereby see different depths, different features and even different phenomena. This is achieved by using different radiations originating in different layers of the solar atmosphere.

2. The outer layers

2.1 *Solar corona*

The outer most layers of the Sun, called the ‘corona’ could be seen only with the help of the Moon when it covers the disk of the Sun during a total solar eclipse (Fig. 3). The long streamers and plumes seen in these pictures are of extremely low density (10^8 /particles per cc) plasma, but the temperature of these features is very high – of the order of a million degrees. The shape and size of the corona changes with time and exhibits solar cycle variations. Figure 3 also shows the variation of coronal structure during maximum and minimum period of the solar cycle. Note the roundish shape observed during the maximum period while during the minimum, the coronal streamers extend more in the east–west direction.

During totality, the naked-eye observers have seen the corona extend to even 15 to 20 solar radii. There is no photographic record available showing such extended corona. This is because the human eye has a very large dynamic range to perceive intensity variations. No photographic or CCD detectors can match it. So to enjoy the beauty of the total solar eclipse, our advice is to just close your eyes, say 20–25 minutes before the totality and get dark-adapted, and open your eyes only at the time of totality. Then you will really see the beauty of the corona. In Fig. 4 is shown a drawing made of the extended corona observed during the 1878 total solar eclipse.

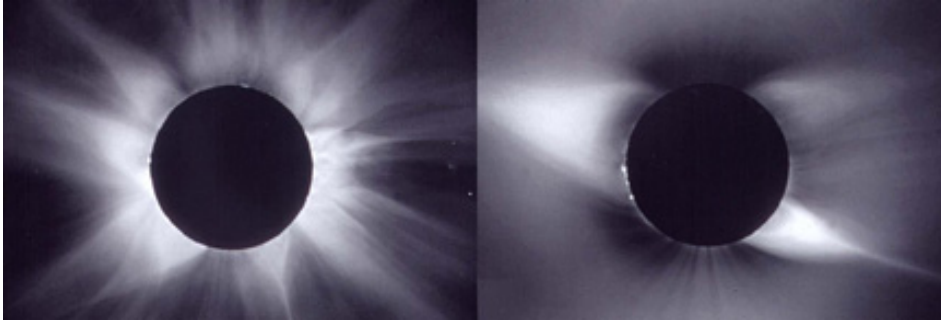


Figure 3. Left picture showing photograph taken on 16 February 1980 during maximum and the right picture taken on 3 November 1994 during minimum activity period.

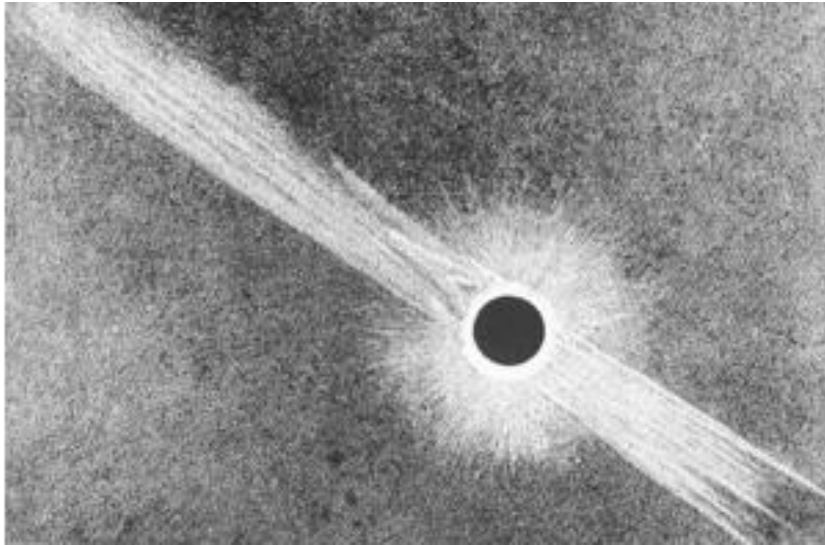


Figure 4. Drawing of the total eclipse made during 27 July 1878, note the extended corona to more than 12–13 solar radii.

2.2 Chromosphere

Using certain radiations, which originate in a particular layer of the solar atmosphere, one can see or detect features and physical conditions of that layer. For example, if you use an instrument to see through the monochromatic light of hydrogen H-alpha or ionized calcium or the radiation from any other strong line formed in the temperature range of about 10,000–12,000 K, you can study the chromosphere. Beyond the chromosphere, the temperature rises very steeply to million degrees in the corona, where most of the atoms get ionized. Radiations originating in the corona or in the transition region (a thin layer between the chromosphere and the corona), are due to highly ionized atoms of iron, calcium, neon, silica, etc., and emit in EUV and X-rays, displaying fine structures in this region of the solar atmosphere. In Fig. 5 is shown full disk pictures taken in the chromosphere in the light of HeI 10,830 Å, EUV at



Figure 5. Full disk pictures taken on the same day in HeI 10830 Å, EUV and Soft X-ray radiations. Note the variation of the disk features in the 3 wavelengths.

195 Å and soft X-ray radiations. Compare the solar structures seen in these 3 wavelengths. In the first image, the large white patch marks the coronal hole region and the dark close structures are where HeI shows absorption. In the X-ray and EUV pictures the bright emission areas indicate close magnetic active regions, with strong magnetic fields.

2.3 Solar prominences

Solar prominences and mass ejections have been observed from early times during solar eclipses. Lately their significance has been realized in view of the Sun–Earth connection. Prominences seen against the disk are called as *filaments*. Extensive study of filaments is being carried out at the non-profit corporation; Helio Research by Sara Martin and her colleagues and using a 10-inch Martin telescope designed for filament observations. In Fig. 6 is shown a huge prominence observed at Big Bear Solar Observatory. Prominences are also observed in EUV radiations and microwave radio spectrum. To really enjoy the dynamical phenomena of solar prominences, one should see the prominence movies now available from several observatories. What supports the prominences, which are at a relatively lower temperature of about 10,000 K, but stay imbedded in the hot corona of million degrees? Why do they remain in the corona for days to weeks and suddenly erupt?

2.4 Coronal mass ejection

Frequently huge volume of plasma of low density are ejected away from the Sun, these are known as Coronal Mass Ejections or CMEs. The first indications of CME came from the Skylab Space mission in the early seventies. A CME seen in action is among the most beautiful dynamic solar phenomenon to observe. Daily movies of CMEs are now taken by LASCO instrument on board the SOHO spacecraft. There are three coronagraphs which can observe CMEs even up to 30 solar radii from the Sun. In Fig. 7 is shown a picture of a huge CME taken from LASCO. How these CMEs are formed and triggered are major problems to solve. Because CMEs are known to generate geomagnetic storms on the Earth, their properties are especially significant to record and study.

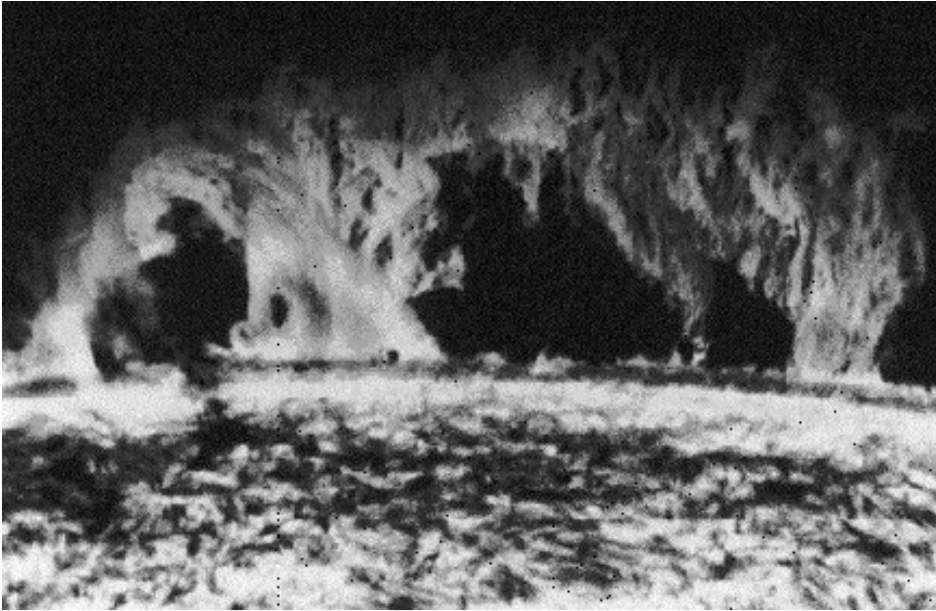


Figure 6. Picture of a huge solar prominence observed at Big Bear Solar Observatory.

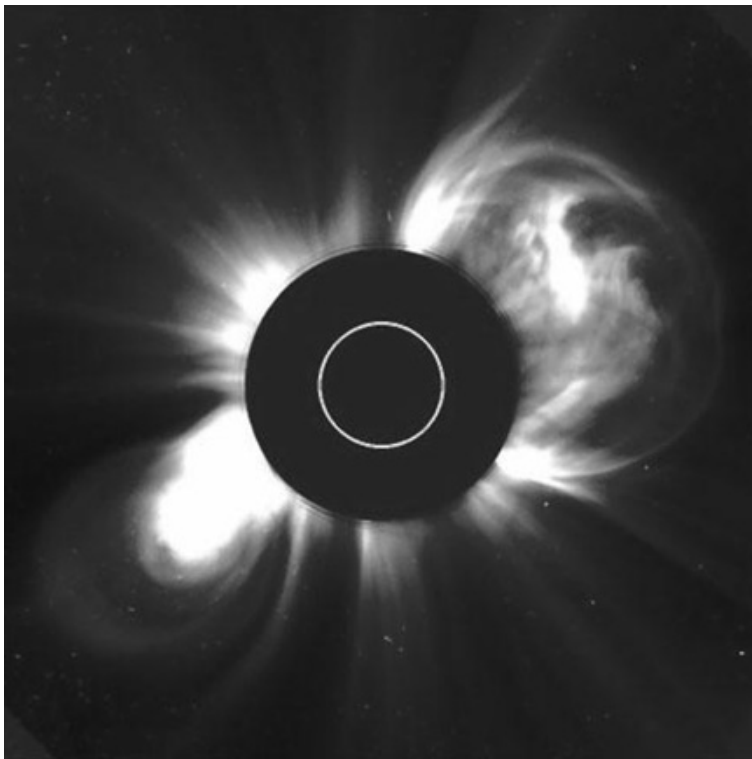


Figure 7. A huge coronal mass ejection observed from LASCO instrument on SOHO spacecraft.

2.5 Coronal holes

The foremost discovery that was revealed by rockets, Orbiting Solar Observatory (OSO) and Skylab spacecrafts in the early seventies was that of large, dark regions observed on X-ray pictures of the Sun, now known as *coronal holes*. Why the X-ray emission is depleted over certain regions on the Sun was a subject of great theoretical investigation. It was further found that these coronal holes correspond to regions where high-speed solar wind emanates from the Sun, and that there is a positional correspondence with the photospheric magnetic field. These regions show unipolar magnetic field, that is the field lines are 'open' over the coronal holes. However, over magnetic active regions the field lines join fields of opposite polarities and these regions appear as intense bright areas in X-rays and EUV. In Fig. 8 is shown a huge coronal hole observed during successive 5 rotations. Notice the rigid rotation of this coronal hole. Why does the coronal hole rotate so rigidly, displaying no differential rotation and appearing almost glued to the photosphere?

3. Photosphere

3.1 Granulation

The whole solar disk appears mottled or granular when seen even under moderate seeing conditions, why? Scientists now believe that it is due to the manifestation of convection from below the photosphere. Although granular structure was observed and beautiful drawings were made by early solar astronomers, its real significance was realized by Schwarzschild and Richardson only in the early fifties. Since then extremely high spatial resolution observations of granulations have been obtained through modern telescopes, such as the Swedish Solar Vacuum Telescope (SSVT),

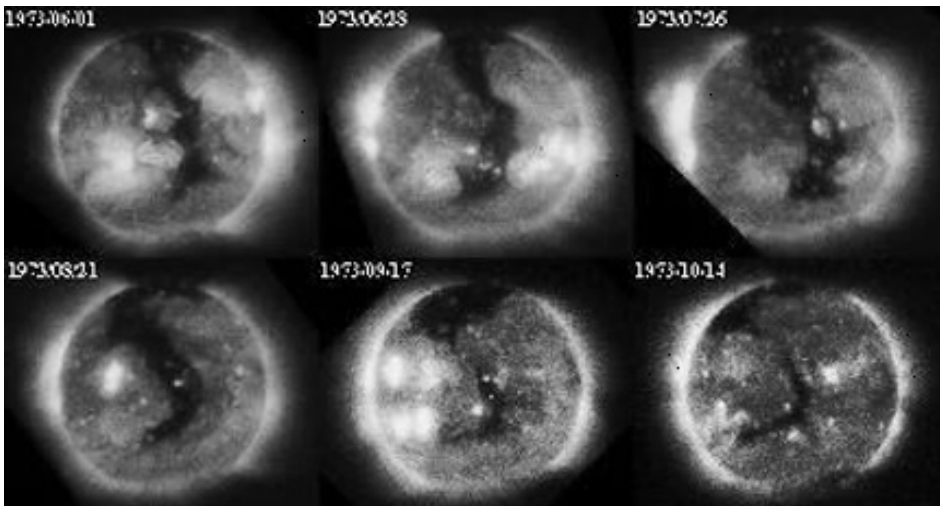


Figure 8. Showing North–South oriented coronal hole on X-ray pictures taken from Skylab, beginning from 1 May 1973 at an interval of 27 days, displaying almost persistent co-rotating coronal hole for 5 rotations.

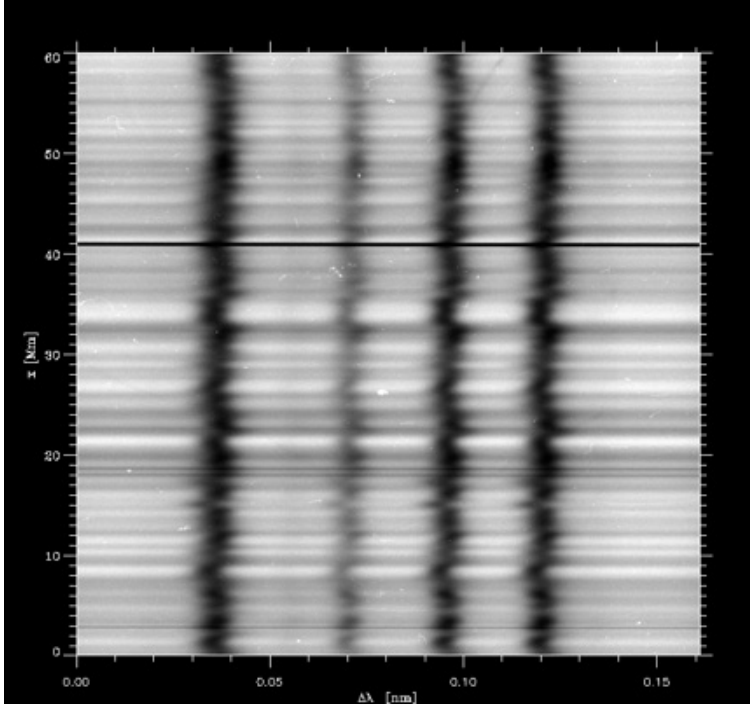


Figure 9. Showing a high resolution spectrum of solar granulation, bright and dark streaks are due to granule brightness variation and zigzag pattern in the absorption lines is due to the vertical granular motion. It will be noticed that the line shifts in darker regions generally appear towards longer wavelength side (right side) implying downward motion.

Dutch Open Telescope (DOT), the German Vacuum Telescope, the Dunn Vacuum telescope at Sacramento Peak in USA and other telescopes. In the left panel of Fig. 1 is shown a high resolution picture taken with the Sac peak telescope. Compare this with Father Secchi's drawing of granulations made in 1878 shown in the right panel of the same figure. How similar they are! The solar granulation also displays vertical motion, which is seen as wiggly lines in the spectrum, due to the Doppler shift of spectral lines as shown in Fig. 9 taken through the German vacuum telescope.

3.2 Sunspots

What more do we see on the solar surface? When one looks in the integrated light, what one generally finds are dark spots of varying size and shapes, lasting from a few hours to several weeks and months. In Fig. 10 is shown a full disk photoheliogram with a large number of spots distributed over the solar disk. Sometimes the Sun is devoid of such dark spots. What are these spots? As they appear dark, they should be relatively cooler regions of the solar atmosphere. They are in fact about 2000 K degrees less than the surrounding photosphere. The questions, are 'how could they stay for weeks and months surrounded by high temperature region and why are they dark?'. It is observed that they are also the seat of strong magnetic field of the order of 2000 to 4000 gauss. Magnetic field in spots was first discovered by George Hale

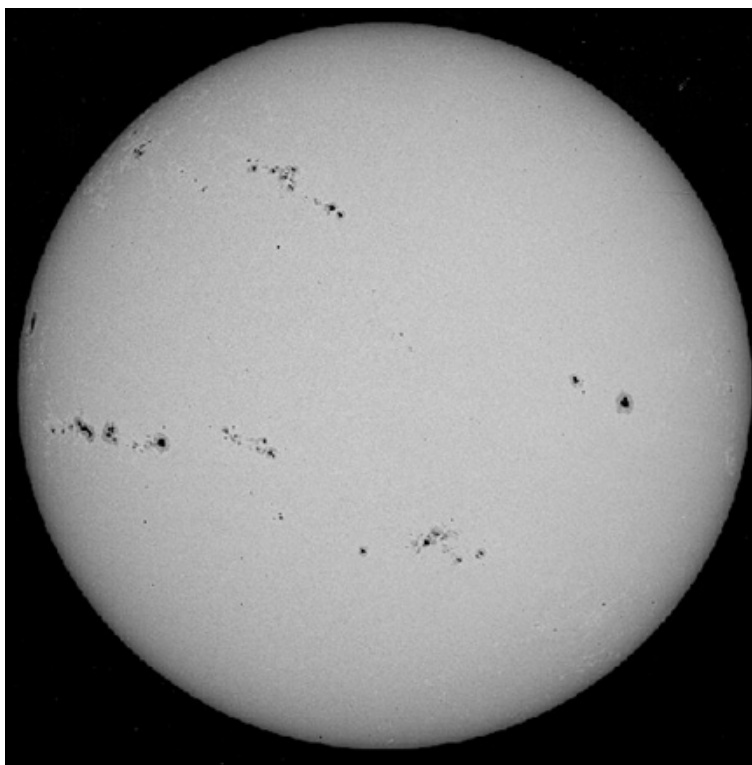


Figure 10. Full disk picture taken in white light, showing large number of sunspots.

at Mount Wilson Observatory in 1907; actually it was the first cosmic magnetic field that was discovered. Some scientists believe that spots are dark because they inhibit the emergence of convective energy due to the magnetic field. This raises the question, ‘what happens to the missing energy from the spots?’.

Some large spots near or on the limb, occasionally display a slight depression that is known as the Wilson Effect. However, it is not yet confirmed whether this is really a physical depression of the solar surface over spots or it is due to a density change. In Fig. 2 is shown a high-resolution picture of a sunspot taken with the Dunn telescope at Sac Peak Observatory. In the same figure is shown Father Secchi’s drawing of a sunspot. Compare the two pictures. How similar they are – one taken with the most modern vacuum telescope and the other drawn by a very skilled observer using an ordinary 13-inch refractor. Sunspots also display intricate structure, both in the penumbrae and umbrae and mass motion. In Fig. 11 is shown a picture of a spot group, taken with the Dunn telescope, displaying beautifully light bridges in the umbra and penumbral filament structures of the order of 0.1 arcsec.

Spot groups appear in different shapes and sizes, in Fig. 12 is shown a peculiar spiral spot group. Such spots develop strong shear and possible evidence of twisting of magnetic field lines. High-resolution movies of sunspots made by DOT and SSVT are now available at DOT and SSVT’s websites on Internet. Some of these movies display penumbral filament motion – inwards towards the umbra. However the Evershed effect is known to display an outward motion. Why are these two oppositely directed motions

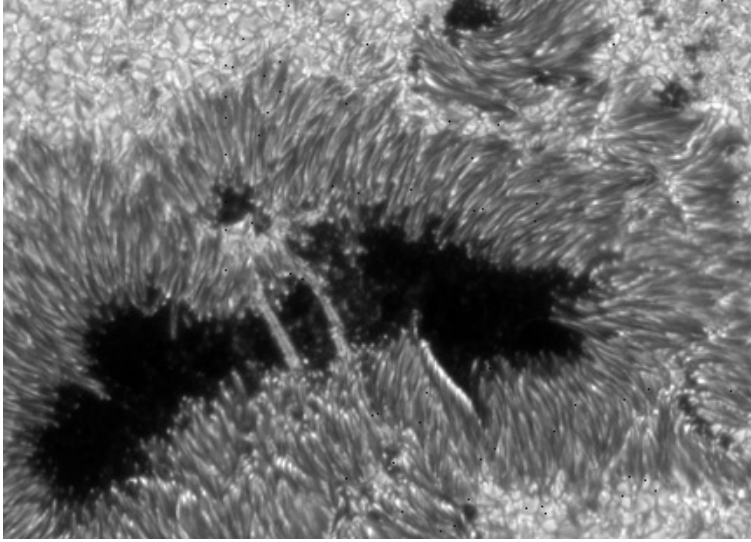


Figure 11. High resolution picture taken from Dunn telescope at Sac Peak of a sunspot group showing light bridges in sunspot umbra.

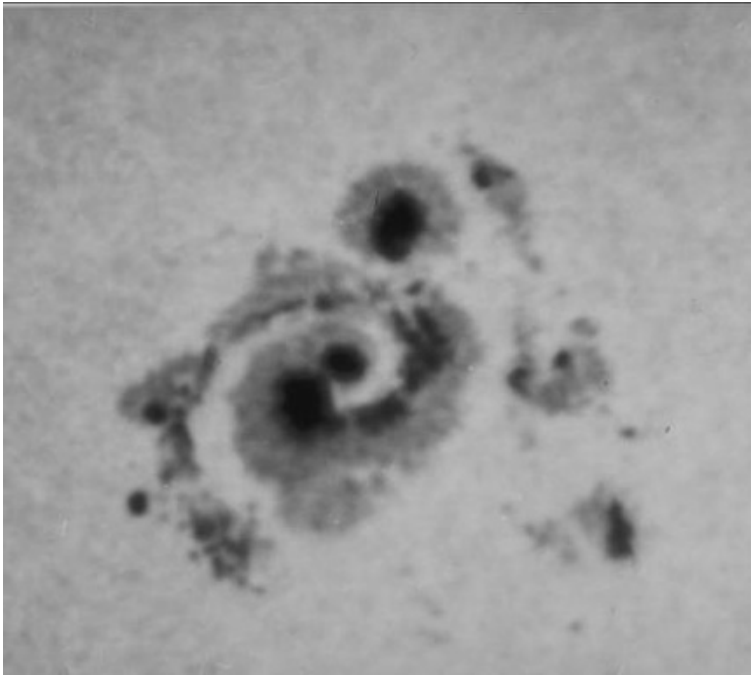


Figure 12. A very peculiar spiral shape sunspot group observed in white light.

existing in the sunspot penumbra? As mentioned earlier sunspots also have magnetic fields, very strong magnetic fields of the order of even 4000 gauss have been detected in spots.

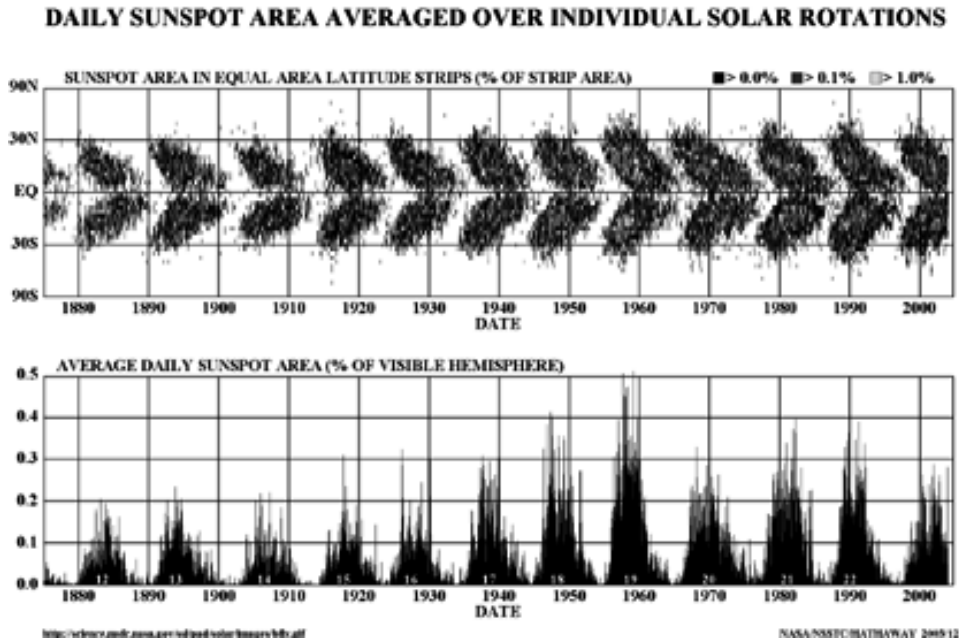


Figure 13. Upper panel shows the ‘Butterfly diagram’ indicating the latitude drift of sunspots during the 11-cycle from 1875 to 2003. Lower panel shows the average daily sunspot numbers during this period.

3.3 Solar cycle

The number of sunspots varies with time, this is known as the *sunspot solar cycle*. Every 11 years our Sun goes through a period with a maximum and minimum number of spots. In Fig. 13 is shown a plot of sunspot position on the disk and number of spots with time. New spots appear at high latitudes and slowly form closer and closer to the solar equator as the cycle progresses. This diagram is known as the ‘butterfly diagram’. Surprisingly during the period from 1640 to 1715 AD, the Sun displayed no solar activity, it was completely devoid of sunspots! This period is known as ‘Maunder Minimum’. Why did the Sun stop producing sunspots during this period? This aspect needs to be investigated, as no explanation has come forward for this minimum. How the sunspots appear and disappear on the solar surface needs a thorough explanation. There are some theories to explain generation and disappearance of sunspots, but none of them are able to explain all the observed phenomena.

3.4 General magnetic field

Besides the strong magnetic field discovered in sunspots, the father and son team, Hecce and Harold Babcocks found that magnetic fields exist all over the Sun, this is now known as the *general magnetic field*. In Fig. 14 is shown a magnetogram of longitudinal – a line of sight component of the magnetic field along with white light picture of that day. The dark areas mark the positive polarity, while the white regions mark the negative polarity. The general magnetic field also shows variation with the solar activity cycle. During high activity periods, the disk is cluttered with large number of

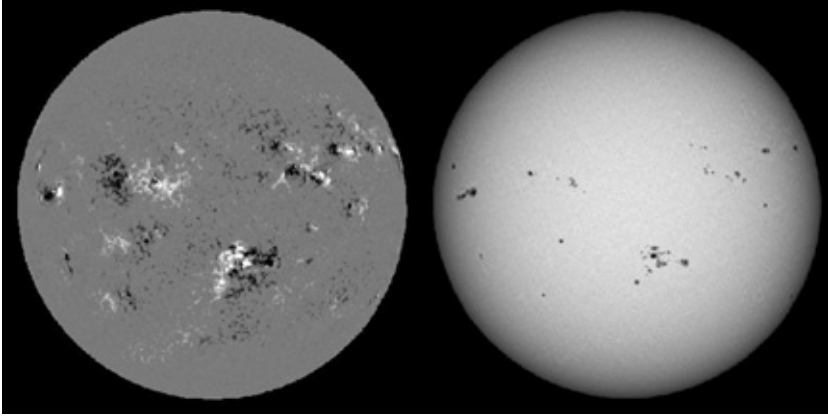


Figure 14. Pictures of full disk white light and magnetogram made during the period of high activity.

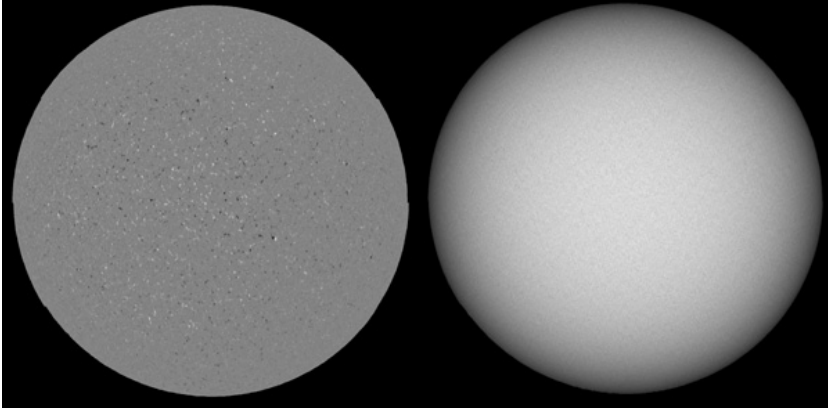


Figure 15. Pictures of full disk white light and magnetogram made during the period of low activity.

sunspot groups and intense magnetic regions, as shown in Fig. 14. During low activity periods, the solar disk appears completely blank without spots. Magnetic fields, as small bipolar mottled structure, as shown in Fig. 15. Maps of the general magnetic field are routinely made and are available on regular basis from number of solar observatories. The GONG (Global Oscillation Network Group) and MDI (Michelson Doppler Imager) instruments provide numerous daily full disk magnetograms. High spatial resolution magnetic field maps are now obtained with SSVT (Swedish Solar Vacuum Telescope) and other modern telescopes displaying magnetic features of less than 0.5 arcsec in sizes. The question of generation and dissipation of magnetic fields on the Sun is one of the puzzling problems for MHD theoreticians.

3.5 Solar irradiance

If the Sun were a variable star, of say only 0.1–0.2-magnitude variability, like many stars in the sky, could the human race have survived? The answer is NO. The solar

radiation has been constant over billions of years, and this has helped to develop life on this planet. Lately, the interest to look into the ‘constancy’ of the *solar irradiance* has greatly increased along with greater interest in Sun–Earth relations. From space-based instruments, highly accurate data are now available which indicate that the solar irradiance, which previously was called as the *solar constant*, is indeed not at all that constant. Slight variation of about 0.1% during the 11-year solar cycle have been noticed. Does this variation affect our Earth’s climate? This is a question to resolve.

4. Solar flares

Occasionally near magnetic active regions, suddenly intense emission of electromagnetic radiations are observed; this phenomenon is known as a *solar flare*. Solar flares are one of the most fascinating phenomena on the Sun. In visible radiations, the best view is in the chromosphere. Solar flares release enormous amounts of energy ranging from 10^{29} to 10^{32} ergs in a few minutes, in spectral frequencies from radio, visible, extreme ultra violet, X-rays and Gamma rays and even particles and cosmic rays. Flares affect our interplanetary space and a whole gamut of our Earth’s environment. One such intense flare was observed at the Udaipur Solar Observatory on 28 October 2003, the H-alpha sequence of development of this double ribbon flare is shown in Fig. 16. Beautiful post flare loops are seen in visible H-alpha and EUV radiations. Extensive study of flare phenomena has been and is still continuing. Both theoretical and experiment minds are engaged in this business. In Fig. 17 is shown a peculiar flare observed by TRACE in EUV; notice the canopy structure of EUV emission.

5. Supergranulation

In the late fifties, a most stunning discovery was made by a physicist, Robert Leighton of Caltech, USA. He demonstrated that over the solar surface there are large ‘velocity cells’ of 30,000–40,000 km size, wherein material flows from the cell’s centre outwards, towards the edge of the cell. These cells could be seen only when one makes a picture of the solar disk displaying Doppler shift of gases. Corresponding to these cells there was nothing seen on pictures taken in white light. However, these cells did correspond in position and structure with network seen in Ca II spectroheliograms. Since then these are called as *supergranulation*. In Fig. 18 is shown a full disk velocity picture displaying supergranulation; the bright and dark regions at the cell boundaries, indicate approaching and receding velocities.

6. EUV emission

Extreme Ultra Violet (EUV) observations made by a number of spacecrafts, namely Skylab, Yohkoh, EIT, TRACE, etc., have revolutionized the observational solar physics. Hundreds of new phenomena have been discovered and studied in great detail. The extremely fine structures of the order of few hundred km in size have been observed by the TRACE, EIT space crafts. These observations reveal that the so called uniform, homogenous solar atmosphere, assumed by theoreticians for simplicity is **not** at all homogeneous, but is permeated by fine structure of loops, controlled by magnetic field. Figure 19 shows a typical example of fine coronal loop structure of less than 0.2 arcsec width. TRACE movies of such loops show the development and dynamics of loops.

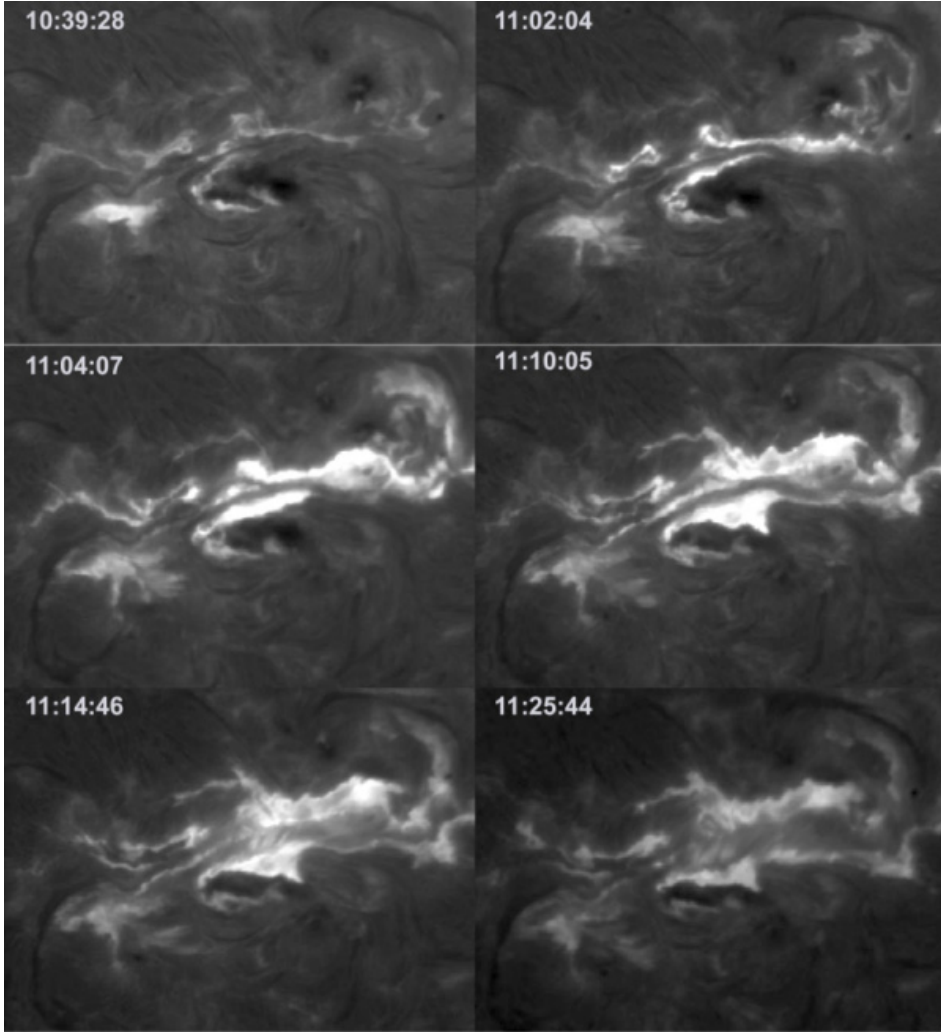


Figure 16. A class 4B flare observed in H-alpha from Udaipur Solar Observatory on 28 October 2003.

7. Helioseismology

Until now we have been discussing and seeing views of the upper and outer layers of the solar atmosphere. Now we will have a glimpse of the solar interior. During the last two decades a new technique has emerged to probe the solar interior called *helioseismology*. This branch of science is the offshoot of Leighton's discovery in 1960 that certain areas or regions on the Sun oscillate with a period of about 5 minutes. These regions may oscillate for about 5 to 6 cycles. Then the oscillations die out but during this time at some other place on the Sun, new sequences of oscillations start. It was realized by Roger Ulrich and by John Leibacher and Robert Stein (Leibacher & Stein 1970) that these observed oscillations are in fact due to the standing pressure p-waves generated inside the Sun. During the last two decades, helioseismology has unraveled several

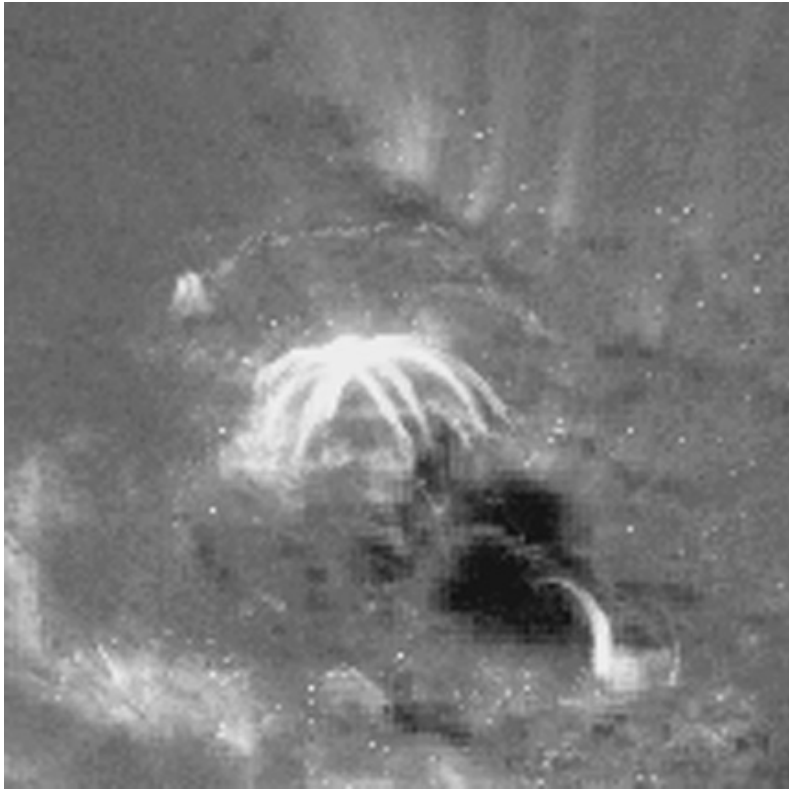


Figure 17. A peculiar flare observed from TRACE on March 17, 2000 showing reconnecting loops.

secrets of our solar interior, solving many puzzling questions like the problem of solar neutrino, Standard Solar Model (SSM), solar rotation, etc. The improved SSM has been put on firm footing through high precision helioseismic observations, and now it does not depend on theoretical assumptions. In Fig. 20 is shown modes of oscillations. Nodes and anti-nodes are shown as dark and bright patches on the solar disk. The cut away portion gives a presentation of p-waves traveling inside the Sun that are reflected at the surface boundary of the Sun. There are more than 10,000 modes of oscillations in the Sun.

The knowledge of rotation of the solar interior is of fundamental importance. Through helioseismic techniques it is now known with a high degree of precision. From the surface down to $0.6R_o$, the Sun shows differential rotation but in the deeper layers it displays solid body rotation. In Fig. 21 is shown the density and sound speed variation with depth from the surface to the deep interior. It is found that around $0.7R_o$ there is 'bump' in the sound speed and density profiles, this region is called the *tachocline*. Why such a bump appears is of great theoretical interest. What does it represent for the internal structure and dynamics of the Sun?

Streams of fast and slow material flows have been observed on the solar surface by Howard and La Bonte in 1980 (Howard & LaBonte 1980) these are called as *torsional oscillations*. In Fig. 22 fast and slow streams are revealed as dark and bright bands.

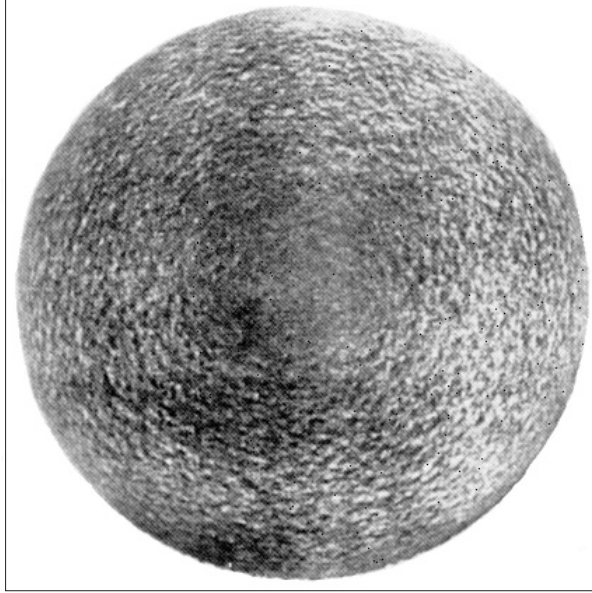


Figure 18. Full disk velocity picture of the Sun displaying supergranulation as cellular structures with dark and white boundaries; away from the central portion the cells are more conspicuous, because the Sun is a spherical body, near the disk centre the cells appear washed out as the motion in the cells is horizontal outwards on the solar surface.

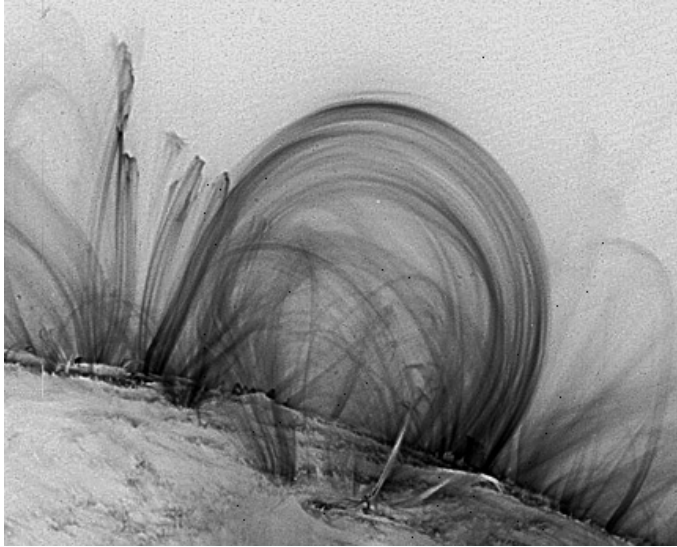


Figure 19. Extremely fine structure coronal arcade of loops observed by TRACE on November 6, 1999.

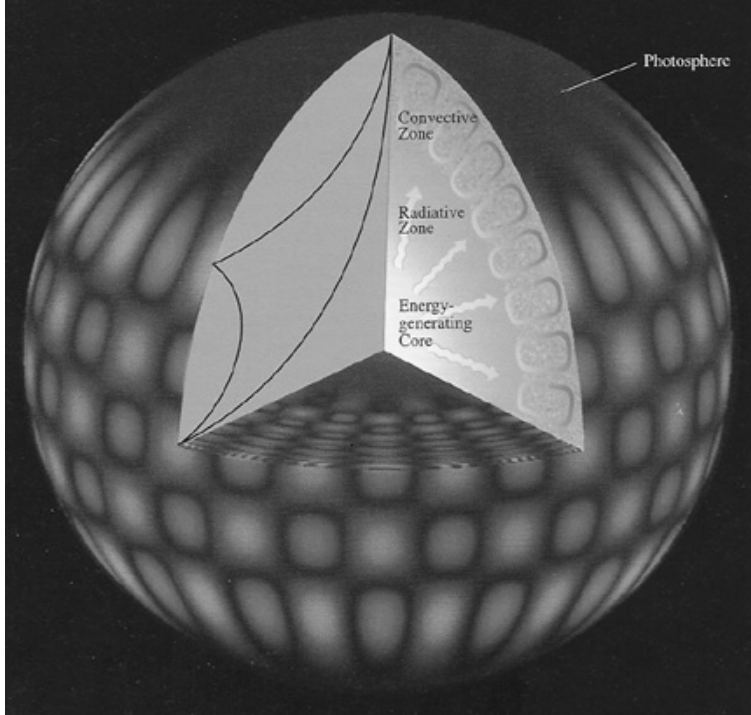


Figure 20. A cutout view of the Sun showing modes of velocity oscillations as dark and gray areas on the solar surface and p-wave traveling inside the Sun, being reflected at the solar surface.

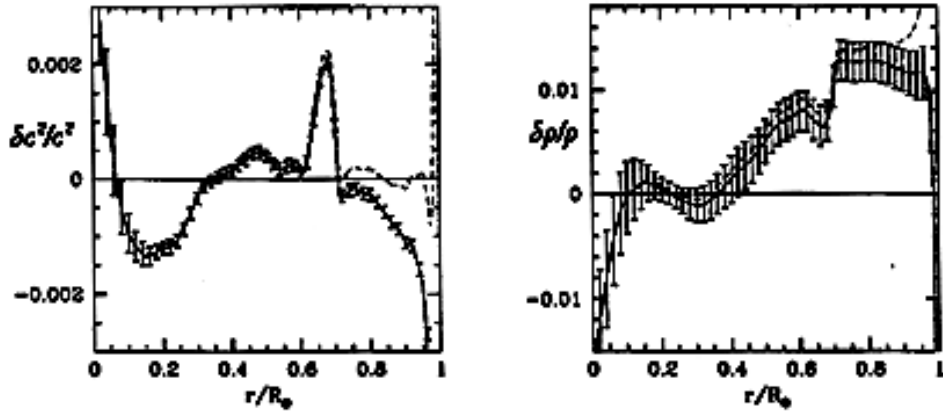


Figure 21. Showing variation of sound speed and density with depth in the solar interior. Note a 'bump' at $0.7R_\odot$, in sound speed and discontinuity in the density profiles, this region is called as 'tachocline'.

These streams move towards the solar equator with advancement of the solar cycle. These streams have been also detected through helioseismic technique, and found to continue in the solar interior at least down to $0.85R_\odot$.

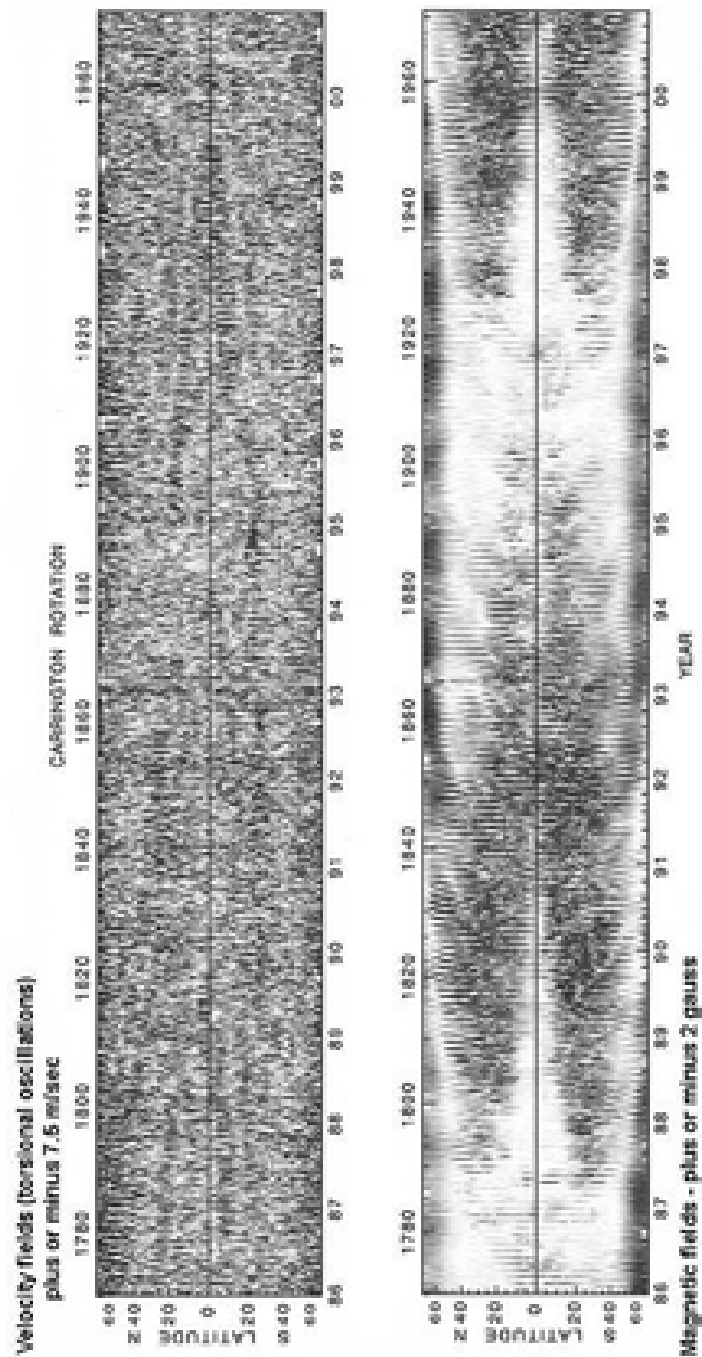


Figure 22. Upper panel shows the torsional oscillations observed on the solar surface during 1986 to 2001 as bright and dark belts, and the lower panel indicates synoptic magnetic field map.



Figure 23. Sun quake observed from MDI on SOHO spacecraft due to a solar flare on 9 July, 1999, note the circular wave-like structure emanating from the flare region.

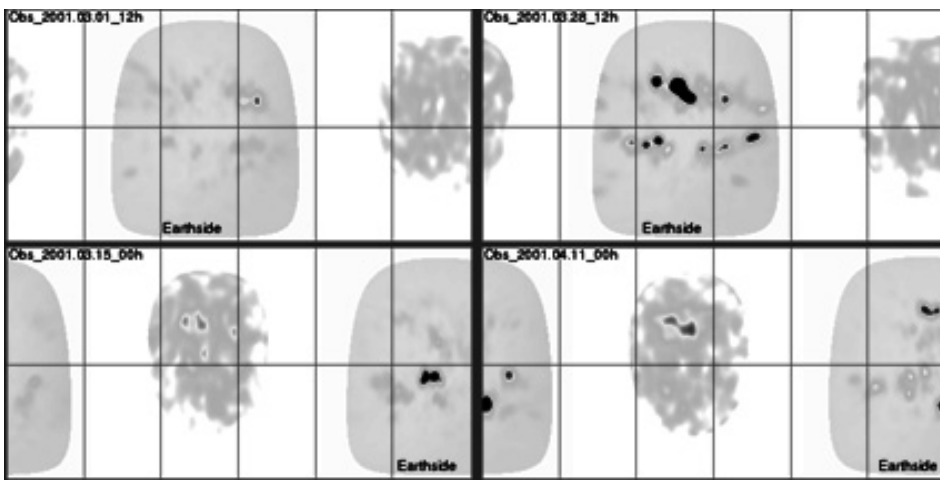


Figure 24. Showing location and structure of active regions on four sample images of the Sun taken from 1 March to 11 April 2001, over a span of 2 rotations. These computed Carrington maps show the magnetic flux on the earth side and inferred flux on the far side of the Sun, was obtained by helioseismic holographic technique at 13.5 days interval.

The effect of flares on solar oscillations have been recently observed in the case of a moderate flare on July 9, 1996 from the MDI instrument on the SOHO spacecraft by the Stanford solar group. This flare produced a ‘wave’ like pattern on the solar surface around the flare region, as shown in Fig. 23. This phenomenon is now known as a ‘Sun quake’. Surprisingly only one such event has been reported in the literature. Helioseismic techniques have helped even to map the ‘other side’ of the Sun – or the side opposite to the front side facing us. In Fig. 24 is shown a view of both the front and backside of the Sun, as it made 2 rotations. These observations are of enormous importance for Space–Weather forecast and in prediction of solar activity, in interplanetary space.

After seeing all these fascinating pictures of our nearest star – the Sun, and knowing that there are still large numbers of challenging and puzzling secrets to unravel, you should be convinced that our Sun is an object in the sky that needs very dedicated, systematic and persistent observations and study.

For this presentation I have heavily drawn from our recent book, 'Fundamentals of Solar Astronomy' (by Arvind Bhatnagar & William Livingston 2005, Publ. by World Scientific Publishing Company, Singapore), therefore I have not given specific references, as they are mentioned in this book.

Acknowledgements

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Some Recent Developments in Solar Dynamo Theory

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Abstract. We discuss the current status of solar dynamo theory and describe the dynamo model developed by our group. The toroidal magnetic field is generated in the tachocline by the strong differential rotation and rises to the solar surface due to magnetic buoyancy to create active regions. The decay of these active regions at the surface gives rise to the poloidal magnetic field by the Babcock–Leighton mechanism. This poloidal field is advected by the meridional circulation first to high latitudes and then down below to the tachocline. Dynamo models based on these ideas match different aspects of observational data reasonably well.

Key words. Sun: magnetic field—sunspots—solar cycle—MHD.

1. Introduction

It appears that major breakthroughs in our understanding of the solar cycle have been taking place approximately at intervals of a half century. The 11-year cycle of sunspots was discovered a little more than 150 years ago (Schwabe 1844); the magnetic nature of this solar cycle became apparent about 100 years ago (Hale 1908); and the hydro-magnetic dynamo theory to explain the origin of this cycle was first formulated about 50 years ago (Parker 1955a). On extrapolating this trend, we expect that another major breakthrough in this field should be taking place right now! We indeed believe this to be the case. Some of the earlier breakthroughs were achieved single-handedly by extraordinary individuals like Hale and Parker. We now live in a less heroic age, and the present breakthrough is due to the joint effort of many groups around the world, in which our group in Bangalore also has played a role. This article is something intermediate between a balanced review article and an exclusive report of the research activities of our group. We shall try to present our work in a historical context against the broader developments in this field. For more elementary and detailed introductions to solar dynamo theory, the readers are urged to look at Choudhuri (1998; Chapter 16) and Choudhuri (2003a).

The basic idea of dynamo theory is that the toroidal and poloidal components of the solar magnetic field give rise to each other in a feedback loop. Since the toroidal component gives rise to the sunspots, we can regard sunspots as a kind of proxy for the toroidal component. On the other hand, the weak magnetic field outside active regions is regarded as a manifestation of the poloidal component. In contrast to the sunspots which migrate equatorward with the solar cycle, this weak field migrates poleward (see, for example, Wang *et al.* 1989). This is clearly seen in Fig. 1, which gives

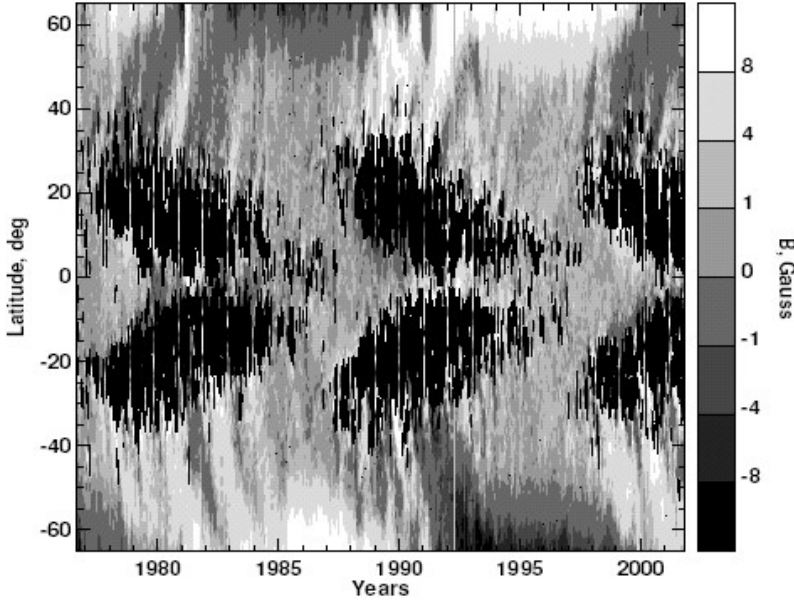


Figure 1. The butterfly diagram of sunspots superposed on a time–latitude plot of the longitude-averaged weak field at the solar surface. Different shades of grey indicate different values of this weak field. Courtesy: A. G. Kosovichev.

time–latitude plot of the longitude-averaged weak field, with the butterfly diagram of sunspots superposed on it. Early work on solar dynamo was mainly confined to explaining the butterfly diagram of sunspots. Now we believe that the behaviour of the weak field also provides some vital clues about the nature of the solar cycle and a full dynamo model should explain this as well.

The generation and the dynamics of the toroidal component is discussed in section 2. Then section 3 discusses the generation of the poloidal component and the role of the meridional circulation. We describe some of the details of our dynamo model (one of the most comprehensive solar dynamo models available at the present time) in section 4. Then our conclusions are summarized in section 5.

2. The generation and dynamics of the toroidal magnetic field

Helioseismology has mapped the angular velocity distribution in the interior of the Sun. A strong radial shear in angular velocity is found at the bottom of the solar convection zone (hereafter SCZ), the region of shear being known as the tachocline. Any magnetic field present in the region of tachocline would be stretched by the differential rotation there to generate a strong component in the toroidal direction. All dynamo theorists agree that the toroidal magnetic field is produced in the tachocline at the base of SCZ. Presumably, interactions with convection keep the magnetic field intermittent, which probably exists in the form of concentrated flux tubes roughly aligned in the toroidal direction (Choudhuri 2003b).

A flux tube in SCZ is subject to magnetic buoyancy (Parker 1955b). However, magnetic buoyancy is significantly suppressed if the flux tube is in a stable region of

subadiabatic gradient below the bottom of SCZ (Parker 1979, section 8.8). Consider a flux tube of which the major part lies in the stable region, with a small portion coming inside SCZ. Then this portion should rise due to magnetic buoyancy to reach the solar surface and should eventually produce a bipolar active region. Such a flux rise can be studied through a simulation of the thin flux tube equation. Since Choudhuri & Gilman (1987) showed that the Coriolis force is extremely important in this problem, this force has to be included in such simulations. Results from such simulation have been reported by different groups (Choudhuri 1989; D'Silva & Choudhuri 1993; Fan *et al.* 1993; Caligari *et al.* 1995).

Bipolar active regions on the solar surface typically tend to be inclined with respect to the solar equator. In spite of a large scatter, this tilt is clearly found to increase with latitude – a result known as Joy's law (Hale *et al.* 1919). One requirement of flux rise simulations is that the upper portions of flux tubes reaching the surface should have inclinations in accordance with Joy's law. D'Silva & Choudhuri (1993) discovered that this requirement is met only if the initial value of the magnetic field inside the flux tube is 10^5G at the starting point at the bottom of the SCZ. In addition to providing the first theoretical explanation of Joy's law nearly three-quarters of a century after its discovery, D'Silva & Choudhuri (1993) provided a tight bound on the value of the magnetic field at the base of SCZ – which was confirmed in subsequent simulations and provided a very important constraint on allowable dynamo models.

3. The generation of the poloidal field and the role of meridional circulation

In his classic paper, Parker (1955a) proposed an idea of how the poloidal magnetic field is generated from the toroidal magnetic field. The convective turbulence, which should be of helical nature in a rotating frame of reference, can drag and twist the toroidal field to give a magnetic field component in the poloidal plane. This idea, which was further developed by Steenbeck, Krause & Rädler (1966), is called the α -effect. However, if the toroidal field is as strong as 10^5G at the base of SCZ – a result forced on us by the flux tube simulations (D'Silva & Choudhuri 1993), then the convective turbulence will not be able to twist this magnetic field, and the α -effect should be completely suppressed for such strong fields.

An alternative idea, which is invoked in many present-day dynamo models, was first suggested by Babcock (1961) and Leighton (1969). Since a bipolar active region appears inclined (Joy's law), the two sunspots with opposite magnetic polarities are at slightly different latitudes. When such a bipolar region decays, the two polarities diffuse preferentially at two slightly different latitudes, thereby giving rise to a poloidal field. The Babcock–Leighton process can be phenomenologically modeled by putting a source term of the poloidal field (represented by an α) near the solar surface. When the usual α -effect was found to be inoperative in the solar interior, it was natural to try using the Babcock–Leighton mechanism to complete the dynamo loop. However, the α representing the Babcock–Leighton process has to be positive in the northern hemisphere – which easily follows from the direction of inclination of active regions. Additionally, $\partial\Omega/\partial r$ is positive within the tachocline at low latitudes. The well-known sign rule of the dynamo (see, for example, Choudhuri 1998, section 16.6) suggests that the dynamo wave should propagate poleward in such a situation, in violation of observations.

Choudhuri *et al.* (1995) showed that it is possible to get an equatorward propagation of the dynamo if there is a suitable meridional circulation. Observations suggest a poleward meridional circulation in the upper layers of SCZ (Giles *et al.* 1997). To conserve mass, there must be a return equatorward flow through the bottom layers of SCZ. This flow was found to force an equatorward propagation of the dynamo, provided the advection time scale of the flow was shorter than the diffusion time scale. This new result of Choudhuri *et al.* (1995) allowed a completely new class of dynamo models to be constructed, with the meridional circulation playing a crucial role in them. Since the poleward flow near the surface advects the poloidal field created at the surface towards the pole, such dynamo models have the additional attractive feature of providing a theoretical model of the observed poleward migration of the weak surface field.

4. The full dynamo model

Our aim now should be to use the helioseismically determined differential rotation to construct a dynamo model, in which the poloidal field is produced by the Babcock–Leighton process and the meridional circulation plays an important role. The toroidal field would be produced in the tachocline and then rise to the surface due to magnetic buoyancy. The poloidal field would be produced at the surface, to be advected by meridional circulation first to the pole and then down underneath to the tachocline where it can be stretched. Two-dimensional dynamo models exploring some of these aspects have been constructed by many authors (Choudhuri *et al.* 1995; Durney 1995, 1997; Dikpati & Charbonneau 1999; Küker *et al.* 2001; Nandy & Choudhuri 2001, 2002; Guerrero & Muñoz 2004; Chatterjee *et al.* 2004).

Helioseismology has found the strongest differential rotation at high latitudes inside the tachocline. When such a rotation profile is used in a dynamo model, the strongest toroidal field is produced at high latitudes and erupts there, contradicting the observations. Nandy & Choudhuri (2002) proposed a way out of this difficulty. Since the meridional circulation is driven by the stresses in SCZ, it is usually assumed to be confined within the SCZ. However, just as we expect some convective overshooting into the stable layers, we can expect the meridional flow to penetrate a little bit into the stable layers below the bottom of SCZ. In that case, even if the strong toroidal field is produced in the tachocline at high latitudes, it is advected by this penetrative meridional flow to lower latitudes through stable layers where magnetic buoyancy is suppressed. Only at low latitudes, it is brought inside the SCZ and rises due to magnetic buoyancy. Nandy & Choudhuri (2002) obtained butterfly diagrams in good agreement with observations, provided the meridional circulation penetrates slightly below the bottom of SCZ.

The solar magnetic field appears to be of dipolar nature. Many of the dynamo simulations were done in one hemisphere, with the boundary conditions at the equator forcing a dipolar solution. One important question is whether the dipolar mode would be favoured in full-sphere simulations. Dikpati & Gilman (2001) claimed that a Babcock–Leighton dynamo preferentially excites a quadrupolar mode. However, Chatterjee *et al.* (2004) showed that the dipolar mode is preferred if the diffusivity of the poloidal component is sufficiently high to couple it between the two hemispheres.

Figure 2 is a theoretical butterfly diagram from Chatterjee *et al.* (2004), showing also the contours of B_r at the solar surface. This figure has to be compared with

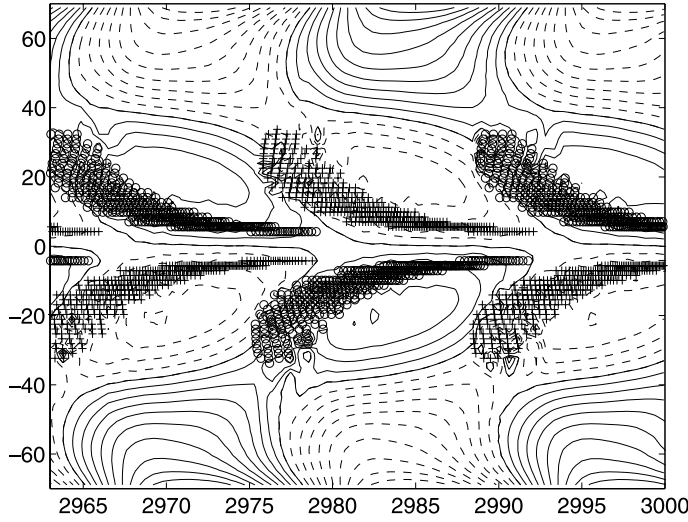


Figure 2. Theoretical butterfly diagram of eruptions from the dynamo model of Chatterjee *et al.* (2004). The background shows contours of the radial field at the solar surface. Eruptions with positive and negative toroidal fields are denoted by '+' and 'o' respectively. The solid contours are for positive B_r and dashed contours for negative B_r .

the observational Fig. 1. It is clear that sunspot eruptions are taking place at correct latitudes. The phase relation between sunspots and the weak field at the surface is also reproduced reasonably well. As we already pointed out, this model favours the dipolar mode.

It may be mentioned that the active regions tend to have negative helicity in the northern hemisphere. One important question is whether this helicity is due to the dynamo and whether our dynamo model can explain it. Based on the qualitative idea of helicity generation proposed by Choudhuri (2003b), it is shown by Choudhuri *et al.* (2004) that their dynamo model also produces the right kind of helicity.

5. Conclusion

In the 1970s and 1980s, many kinematic solar dynamo models were developed by assuming different hypothetical profiles of α and Ω . Now we know that most of these models were based on assumptions which do not correspond to reality. An attractive aspect of the present-day dynamo models is that many of the crucial ingredients of the model are directly based on observations. We know the distribution of Ω from helioseismology. We also observe the decay of tilted bipolar regions on the solar surface and see the Babcock–Leighton mechanism in action. The meridional circulation in the upper layers of SCZ also has been mapped. One of the big remaining uncertainties at the present time is that we do not have any observational results about the nature of the meridional circulation in the lower layers of SCZ. As pointed out by Nandy & Choudhuri (2002), the solutions can be very different depending on whether this circulation penetrates in the stable layers below the bottom of SCZ. This has become an intensely controversial subject at present (Dikpati *et al.* 2005; Choudhuri *et al.* 2005). We hope that future observations and more detailed theoretical modeling will eventually settle

this issue. In spite of many remaining gaps in our understanding, the dynamo models developed in the recent years are the first sufficiently-detailed and realistic models of the solar dynamo, of which many aspects will probably stand the test of time. That is why we regard the recent developments as another breakthrough in the historical evolution of solar dynamo theory, as we mentioned at the very beginning of this article.

Although the traditional α -effect would be quenched at the bottom of SCZ with concentrated magnetic fields of order 10^5 G, there can be other physical mechanisms there responsible for giving rise to effects similar to the α -effect. Some studies are being made of such interface dynamos with α at the bottom. We do not discuss these interface dynamo models in this brief article.

Finally, many aspects of the dynamo models we discussed in section 4 are yet to be explored. Since the magnetic field is expected to be intermittent within the SCZ, one important issue is to understand the relation between the dynamo process and the flux tubes. An initial exploratory study of this problem has been presented by Choudhuri (1993b). Another important problem is to understand the irregularities of the solar cycle – at least at a qualitative level. We are now in the process of incorporating stochastic fluctuations in our dynamo model to study this problem.

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I wish to thank Piyali Chatterjee and Dibyendu Nandy for many useful discussions on the solar dynamo problem.

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Helicity of Solar Active Regions from a Dynamo Model

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Abstract. We calculate helicities of solar active regions based on the idea that poloidal flux lines get wrapped around a toroidal flux tube rising through the convection zone, thereby giving rise to the helicity. We use our solar dynamo model based on the Babcock–Leighton α -effect to study how helicity varies with latitude and time.

Key words. MHD—Sun: activity—Sun: magnetic fields—sunspots.

1. Introduction

Typically, solar active regions are known to have helicity associated with them and the observational studies based on vector magnetograms (Seehafer (1990); Pevtsov *et al.* (1995, 2001); Abramenko *et al.* (1997) and Bao & Zhang (1998)) indicate that the preferred sign of this helicity is opposite in the two hemispheres (negative in northern hemisphere and positive in the southern hemisphere), in spite of a very large statistical scatter. Figure 2 of Canfield & Pevtsov (2000) is a typical plot showing a variation of helicity with latitude, which any theoretical model has to explain.

Solar magnetic fields are believed to be produced by the dynamo process. One possibility is that the dynamo process itself is responsible for the generation of helicity. The other possibility is that the rising flux tubes, which eventually form active regions, get the twist by interacting with the helical turbulence in the surrounding convection christened the Σ -effect by Longcope *et al.* (1998). The two possibilities mentioned above need not be mutually exclusive: both may be simultaneously operative. A careful comparison between observational data and detailed theoretical models will be needed to ascertain the relative importance of these two effects.

We present here calculations of helicity based on our two-dimensional kinematic solar dynamo model presented in Nandy & Choudhuri (2002) and Chatterjee *et al.* (2004).

The dynamo equation deals with the mean magnetic field (see, for example, Choudhuri 1998, Ch. 16), whereas we want to find helicities of active regions which form from flux tubes. To make a connection between these two, we have to look at the relation between dynamo theory and flux tubes. In a Babcock–Leighton dynamo poloidal field A is created from decay of tilted active regions, the amount of tilt being given by the Joy’s Law. Also in accordance with the Hale’s polarity rule, a bipolar active region formed by a flux tube with positive B_ϕ would have the leading spot towards the equator than the following spot. Decay of such a pair would thus mean clockwise lines of B_p around active regions. When a new toroidal flux tube with positive B_ϕ moves upwards

near the surface, the poloidal field gets wrapped around the toroidal flux tube. Due to high magnetic Reynolds number the flux tubes are not able to cut through the poloidal field lines thus giving rise to the helicity.

2. Estimating the value of helicity

For force free fields in the photosphere and the corona we may define helicity as

$$\alpha = \frac{(\nabla \times \mathbf{B})_z}{B_z}, \quad (1)$$

where z corresponds to the vertical direction, which is along the axis of the flux tube for active regions on the surface. The parameter α (not to be confused with the dynamo α -effect traditionally associated with the poloidal field generation mechanism) is a measure of the handedness or chirality of the magnetic field. This twist parameter is an indicator of how stressed the active region flux system is and it is known to play an important role in the flaring and explosive activity of active region magnetic fields (Canfield *et al.* 1999; Nandy *et al.* 2003). The typical observed value of the twist parameter α from magnetogram data, calculated by several authors is about $2 \times 10^{-8} \text{ m}^{-1}$.

To estimate the value of helicity theoretically, we have to keep in mind that the flux of poloidal field B_P through the whole SCZ gets dragged by the toroidal flux tube rising under magnetic buoyancy (see Fig. 4 of Choudhuri 2003). If d is the depth of the convection zone, the flux dragged by the tube is

$$F \approx B_P d. \quad (2)$$

This flux F gets wrapped around the tube of radius a . In an ideal-MHD situation, this flux F would be confined to a narrow sheath around the flux tube. In reality, however, we expect that the turbulence around the flux tube would make this flux F penetrate into the flux tube. Then the magnetic field going around the tube can be taken to be of order F/a . The current density $|\nabla \times \mathbf{B}|$ associated with this field is of order F/a^2 and is along the axis of the tube. If B_T is the magnetic field inside the flux tube, then it follows from (1):

$$\alpha \approx \frac{F/a^2}{B_T} \approx \frac{B_P d}{B_T a^2} \quad (3)$$

on substituting from (2) for F . We use $B_P \approx 1 \text{ G}$, the depth of the SCZ $d \approx 2 \times 10^8 \text{ m}$ and the field inside sunspots $B_T \approx 3000 \text{ G}$. On taking the radius of the sunspot $a \approx 2000 \text{ km}$ and $a \approx 5000 \text{ km}$, we get $\alpha \approx 2 \times 10^{-8} \text{ m}^{-1}$ and $\alpha \approx 3 \times 10^{-9} \text{ m}^{-1}$ respectively. Thus, from very simple arguments, we get the correct order of magnitude.

3. Results from dynamo simulation

In section 4 of Chatterjee *et al.* (2004) we have presented a particular dynamo model which we refer to as our *standard model*. We now present helicity calculations based on this standard model.

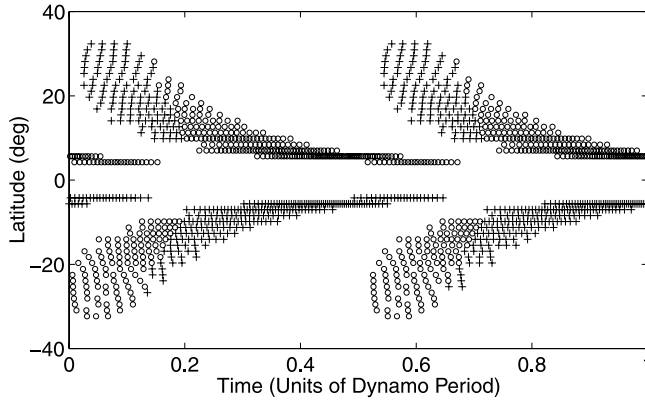


Figure 1. Theoretical butterfly diagram of eruptions from our standard dynamo model. Eruptions with positive and negative helicities are denoted by ‘+’ and ‘o’ respectively.

A flux eruption takes place in our model whenever the toroidal field at the bottom of the SCZ exceeds a critical value. Whenever an eruption takes place in our dynamo simulation, we calculate the poloidal flux F through SCZ at the eruption latitude by integrating B_θ from the bottom of SCZ ($r = R_b$) to the top ($r = R_\odot$), i.e.,

$$F = \int_{R_b}^{R_\odot} B_\theta dr.$$

The value of F calculated at the eruption latitude at the time of eruption gives the amplitude of helicity associated with the eruption. If the sign of F is opposite to the sign of the toroidal field B at the bottom of SCZ, then the helicity is taken as negative (otherwise it is positive). Figure 1 shows the simulated butterfly diagram, indicating active regions of positive and negative helicity. During the solar maximum, the helicity is negative in the northern hemisphere and positive in the southern, as we expect. However, at the beginning of a cycle, there is a short duration when the sign of helicity is ‘wrong’, i.e., opposite of the preferred helicity. We find that 67% of the eruptions in our simulation have ‘correct’ helicity.

Figure 2(a) is a plot of helicity associated with eruptions at different latitudes. This is the theoretical plot that has to be compared with observational plots like Fig. 2 of Canfield & Pevtsov (2000). We notice that the theoretical plot has considerably less scatter compared to the observational data.

To see the variation of helicity with the cycle, Fig. 2(b) and Fig. 2(c) present plots of helicity for eruptions during 4 years of solar maximum and 4 years at the beginning of the cycle respectively. The straight lines represent the least-square fits. For ‘correct’ helicity (negative in north and positive in south), the gradient $d\alpha/d\lambda$ of the straight line has to be negative, as we see in Fig. 2(b) corresponding to solar maximum (λ is the latitude). The gradient, however, is positive at the start of the cycle. To find out how this gradient varies with the cycle, we divide the cycle period into 16 equal intervals and then find the gradient $d\alpha/d\lambda$ for each of the intervals by using eruptions during that interval. Figure 3 shows how the gradient $d\alpha/d\lambda$ varies with the solar cycle. If the Σ -effect makes a significant contribution in the production of helicity, then the variation with the cycle may be less pronounced compared to what we find in our model without

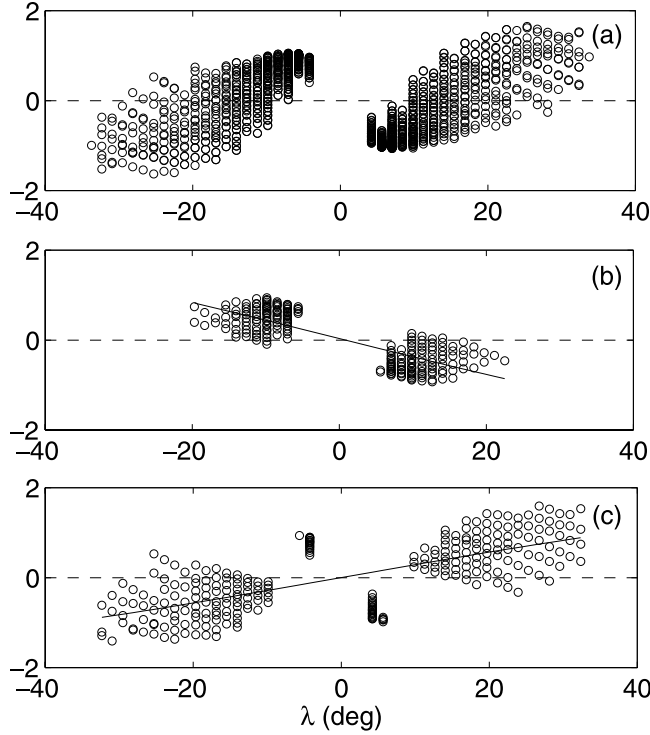


Figure 2. Helicity α (plotted along the vertical axis in arbitrary units) for eruptions at different latitudes, denoted by open circles: **(a)** for an entire solar cycle; **(b)** for 4 years during the maximum and the declining phases of the solar cycle; **(c)** for the first 4 years of the solar cycle. The solid lines in (b) and (c) are least-square fits to the model results.

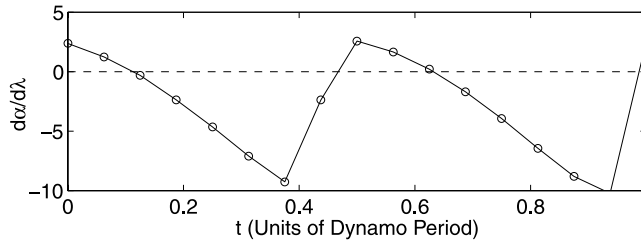


Figure 3. $d\alpha/d\lambda$ as a function of time covering the equivalent of two sunspot cycles. To find out the values of time which correspond to maxima or minima, look at Fig. 1 which has the same horizontal axis.

this effect, since the Σ -effect is cycle-independent (Longcope *et al.* 1998). Since it is in principle possible to determine from observational data how $d\alpha/d\lambda$ actually varies with the solar cycle, a plot like Fig. 3 provides a powerful tool for comparison between theory and observations. There is some indication in the existing data that $d\alpha/d\lambda$ may be varying in accordance with our model; but the data are noisy and the results from different instruments often diverge widely, so it is difficult to draw firm conclusions at this point in time (Pevtsov, private communication).

4. Conclusion

Two clear theoretical predictions follow from our model.

- (1) Since the helicity goes as a^{-2} as seen from (3), the smaller sunspots should statistically have stronger helicity (i.e., higher values of twist α).
- (2) At the beginning of a cycle, helicity should be opposite of what is usually observed.

The alternative model for the generation of helicity, the Σ -effect proposed by Longcope *et al.* (1998), also makes the prediction (1), but not (2). Since helicity is imparted to the flux tubes by helical turbulence of SCZ in this model, the helicity is not expected to vary with the solar cycle. A careful analysis of any possible variation of helicity with the solar cycle would be the best way of ascertaining relative contributions of the Σ -effect and the dynamo (the process studied in this paper) in generating helicity.

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Nanoflares and Heating of the Solar Corona

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Abstract. Coronal heating by nanoflares is presented by using observational, analytical, numerical simulation and statistical results. Numerical simulations show the formation of numerous current sheets if the magnetic field is sheared and bipoles have unequal pole strengths. This fact supports the generation of nanoflares and heating by them. The occurrence frequency of transients such as flares, nano/microflares, on the Sun exhibits a power-law distribution with exponent α varying between 1.4 and 3.3. For nanoflares heating α must be greater than 2. It is likely that the nanoflare heating can be reproduced by dissipating Alfvén waves. Only observations from future space missions such as Solar-B, to be launched in 2006, can shed further light on whether Alfvén waves or nanoflares, heat the solar corona.

Key words. Sun: stars—magnetic reconnection: coronal heating—nanoflares.

1. Introduction

It is well-known that the solar corona is a few hundred times hotter than the underlying atmospheric regions (chromosphere and photosphere). Two heating agents involving magnetic fields, namely MHD waves and transients (such as flares, nano- and microflares) are the most favoured ones (see, e.g., Narain & Ulmschneider 1996; Pandey & Narain 2001; Dwivedi 2002, 2003; Walsh & Ireland 2003; Aschwanden 2004 and references therein). In this article heating by nanoflares (Parker 1988, 1991) is briefly described and discussed.

Section 2 estimates the energy of a nanoflare as a result of a single magnetic reconnection. In section 3, we exhibit the power-law distribution of transients. Section 4 contains a brief description of observational and theoretical efforts made in this direction. The last section contains our conclusions.

Throughout the cgs system of units has been used.

2. Energy of a nanoflare

It is thought that when two oppositely directed magnetic fields come closer to form a current sheet, the current density of contained plasma increases considerably so that a small resistivity is quite sufficient to convert magnetic energy to thermal energy via magnetic reconnection and resulting turbulence. This is shown in Fig. 1.

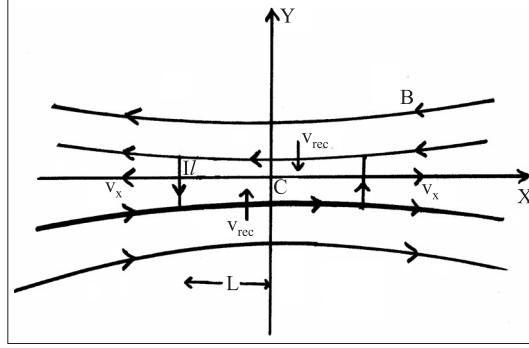


Figure 1.

The current sheet has length $2L$ and breadth $2l$. The magnetic field component B_X reverses direction along the line $y = 0$, assuming field geometry to be independent of coordinate z (directed out of paper) following Tandberg-Hanssen and Emslie (1988), the current density in the current sheet is given by:

$$j = \frac{cB_X}{4\pi l}, \quad (1)$$

where j is the current density and c is the speed of light in vacuum. Equation (1) shows that as l becomes smaller j becomes larger. The merging of oppositely directed field lines, which decreases l , is limited by the gas pressure p between the oppositely directed fields. Since the field and plasma are frozen together, the following equation is satisfied:

$$p = \frac{B_X^2}{8\pi}. \quad (2)$$

Outside the region $|x| > L$, the gas pressure is much smaller than inside the current sheet, $|x| < L$, so that the fluid is ejected along the field lines (x -direction, Fig. 1), reducing the built-up pressure.

In the steady state the momentum equation for the plasma, ejected along the field lines is

$$\rho v_x \left(\frac{\partial v_x}{\partial x} \right) = \frac{-\partial p}{\partial x}, \quad (3)$$

where ρ is the matter density of fluid. Assuming fluid to be incompressible, equation (3) may be integrated from $x = 0$ to $x = L$, corresponding to $p = p_i$ and $p = p_0$, to get

$$0.5\rho v_x^2(L) = p_i - p_0. \quad (4)$$

By symmetry $v_x(x = 0) = 0$ and the pressure difference $p_i - p_0$ is due to the magnetic field pressure $B_x^2/8\pi$ within the region $|x| \leq L$ so that equation (4) gives

$$v_x(L) = \frac{B_x}{(4\pi\rho)^{0.5}} = v_A, \quad (5)$$

where v_A is the Alfvén speed within the reconnecting region. The reconnection velocity $v_{\text{rec}} (= dl/dt)$ gives the rate at which the magnetic field lines are swept into the reconnecting volume, V .

The continuity equation, under incompressibility conditions, demands that the outward flow of matter along the x -axis must be balanced by an inflow in the y -direction so that

$$v_{\text{rec}}L = v_A l. \quad (6)$$

Further, in the steady state, the ohmic dissipation, ηj^2 , of energy in the reconnecting region must be just sufficient to balance the influx of magnetic energy in that region, i.e.,

$$\iiint_V \eta J^2 dV = \iint_S \left(\frac{B_x^2}{8\pi} \right) v_x dS$$

which, in view of equation (6), Fig. 1 and per unit length in the z -direction gives

$$j^2 = \frac{B_x^2 v_{\text{rec}}}{8\pi \eta l}, \quad (7)$$

where η is the resistivity of the fluid in the current sheet. Comparing equations (1) and (7) leads to

$$v_{\text{rec}}^2 = \frac{\eta c^2 v_A}{2\pi L} \quad (8)$$

and

$$l^2 = \frac{\eta c^2 L}{2\pi v_A}. \quad (9)$$

The longitudinal magnetic Reynolds number, R_{ml} , provides the effect of diffusion over dissipation and is defined by

$$R_{ml} = \frac{|\vec{\nabla}x(\vec{v}x\vec{B})|}{|\eta c^2 \nabla^2 \vec{B}/4\pi|} \approx \frac{8\pi L v_A}{\eta c^2}. \quad (10)$$

Equations (8), (9) and (10) now give

$$v_{\text{rec}} = \frac{2v_A}{R_{ml}^{0.5}} \quad (11)$$

and

$$l = \frac{2L}{R_{ml}^{0.5}}. \quad (12)$$

Defining Alfvén crossing time, t_A , and reconnection time, t_R , by

$$t_A = \frac{l}{v_A} \quad (13)$$

and

$$t_R = \frac{l}{v_{\text{rec}}}, \quad (14)$$

equations (11) and (12) give

$$t_R = 0.5 t_A R_{ml}^{0.5}. \quad (15)$$

The rate of energy release in the current sheet may be obtained from the following:

$$\frac{dE}{dt} = \frac{B^2 V}{8\pi t_R}, \quad (16)$$

where B is the magnetic field strength before reconnection (dissipation). The resistivity, η , for the solar corona is determined using the following expression (Spitzer 1962):

$$\eta = 1.5 \times 10^{-7} T^{-1.5}. \quad (17)$$

For a rough estimate, we take $L = 10^9$ cm, $B = 300$ G, electron/proton number density $n = 10^{10}$ cm $^{-3}$ and $T = 2$ MK. This gives:

$$v_A = 6 \times 10^8 \text{ cm s}^{-1}, \quad \eta = 5 \times 10^{-17} \text{ s}, \quad R_{ml} = 10^{14}, \quad l = 200 \text{ cm},$$

$$t_A = 3 \times 10^{-5} \text{ s} \quad \text{and} \quad t_R = 2 \text{ s}.$$

Taking $V = L^2 l = 2 \times 10^{20}$ cm ^{2} , equation (16) together with above values gives:

$$\frac{dE}{dt} = 4 \times 10^{23} \text{ erg s}^{-1}. \quad (18)$$

Thus the energy released from the magnetic reconnection in a single current sheet equals that of a nanoflare.

3. Power-law distribution

Parker (1988) proposed that the energy dissipation of the stressed magnetic structure takes place in a large number of small events, which he termed nanoflares. The superposition of a large number of such events may give the global appearance of a spatially homogenous and stationary heating process. The spiky (both in space and time) heating of nanoflares may be related to global heating as follows:

If E_t be the total heating rate contributed by all the events in the energy range $E_{\min}$, $E_{\max}$ then

$$E_t = \int_{E_{\min}}^{E_{\max}} P(E) E dE, \quad (19)$$

where $P(E)$ is the number of events per unit energy range and time. This occurrence rate displays the following power-law behaviour (Pandey & Narain 2001 and references therein):

$$P(E) = A E^{-\alpha}, \quad (20)$$

where A and α are some constants. On combining equations (19) and (20) we get:

$$E_t = \frac{A(E_{\max}^{2-\alpha} - E_{\min}^{2-\alpha})}{(2-\alpha)}. \quad (21)$$

3.1 Case I: $\alpha < 2$

Since $E_{\max} \gg E_{\min}$, therefore the second term in equation (21) may be dropped to get

$$E_t^I = \frac{A E_{\max}^{2-\alpha}}{(2-\alpha)}. \quad (22)$$

That is, the heating will be dominated by high energy events, namely large and intermediate energy flares.

3.2 Case II: $\alpha > 2$

In this case the first term in equation (21) may be dropped to get

$$E_t^{II} = \frac{A E_{\min}^{2-\alpha}}{(\alpha-2)}. \quad (23)$$

That is, the heating is dominated by the low-energy events, namely micro- and nanoflares.

4. Results and discussion

It seems quite interesting to start with Georgoulis and Vlahos (1996) who developed a cellular automation self-organised critical (SOC) model for transients in which an explosion affects its neighbourhood by lowering the instability criteria and igniting secondary bursts triggered by an initial instability in an avalanche type manner. In the resulting frequency distributions (cf. equation (20)) they obtain two distinct power laws. The weaker events have shorter and steeper power law with exponent $\alpha \approx 3.26$ whereas the large and intermediate events have $\alpha \approx 1.73$. Also the weaker events are responsible for almost 90% of the total magnetic energy released and thus support coronal heating by nanoflares. This model has been further improved by Georgoulis *et al.* (2001).

In order to know whether heating is dominated by low or high energy events we compile the values of power-law exponent, α in Table 1 for the solar corona. It contains both, the theoretically and observationally deduced values.

An examination of Table 1 shows that both types of cases, i.e., $\alpha < 2$ and $\alpha > 2$ exist. No clear picture emerges. There is some controversy over the method of determination of values of α (see, e.g., McIntosh & Charbonneau 2001; Aschwanden & Charbonneau 2002a).

Karpen *et al.* (1996) and Antiochos *et al.* (2002) have performed numerical simulations of the interaction between two bipoles through magnetic reconnection in the lower solar atmosphere. Here the magnetic field is sheared asymmetrically and the biopoles have markedly unequal field strengths. They arrive at the most exciting and potentially far-reaching discovery that the random nature of reconnection process creates a distribution of current sheets throughout the region occupied by reconnected field lines. Reconnection between the larger and smaller flux systems leads to long-lived current sheets that decay slowly and yield the observed X-ray structures.

Table 1. The power-law exponent, α .

Object	Phenomenon/instrumentation	Value	References
Sun	Type I radio burst	3	Mercier & Trottet (1997)
Sun	Iron lines/EIT (SOHO)	2.3–2.6	Krucker & Benz (1998)
XBP (Sun)	X-ray/Yohkoh	1.7 ± 0.4	Shimojo & Shibata (1999)
Quiet Sun	171 Å, 195 Å/TRACE	1.83 ± 0.07	Aschwanden <i>et al.</i> (2000)
Quiet Sun	EUV/CDS (SOHO)	2.5 (Cell) 1.7 (Network)	Harra <i>et al.</i> (2000)
Sun	CIII 977 Å, NIV 765 Å, OVI 1032 Å, NeVIII/SUMER	2.9 ± 0.1	Winebarger <i>et al.</i> (2002)
Solar coronal loop	2d-MHD/simulation	1.5	Dmitruk & Gomez (1997, 1999)
Solar coronal loop	1d-MHD compressible/simulation	1.57 (large events) ≥ 1.81 (nanofalres)	Galtier (1999)
Solar corona	Flarelike processes 0.7–1.1 MK 1.0–1.5 MK 2.0–4.0 MK Hard X-rays	1.86 ± 0.07 1.81 ± 0.10 1.57 ± 0.15 1.4–1.6	Aschwanden & Parnell (2002b)
Solar corona	SOC Model/weak events large and intermediate events	3.26 1.73	Georgoulis & Vlahos (1996)

Another interesting and important result comes from Moriyasu *et al.* (2004) who examine the creation of hot corona in an initially cool loop as a result of nonlinear Alfvén waves. Here the dissipation takes place via mode-coupling with slow- and fast-mode waves which is balanced by conduction and radiative cooling. Their numerical simulations reveal that the resulting corona is very dynamic and full of shocks so that the temporal variation of the X-ray and EUV intensities shows many nanoflare-type events, quite similar to what is actually observed. In fact, the occurrence frequency of X-ray and EUV nanoflares as a function of their peak intensity is a power-law distribution with an index ~ 1.7 for X-ray and ~ 1.39 for EUV. Quite obviously, the actual observed time variation of the X-ray and EUV fluxes in the corona and in the chromosphere, may not be evidence of small scale magnetic reconnections but could actually be due to Alfvén waves.

In another landmark work Cargill and Klimchuk (2004) investigate the radiative signatures of the nanoflare-heated corona and speculate that if an observed coronal loop contains many small strands then continual heating and cooling of strands would lead to a corona having a multi-thermal structure. The cyclical heating-cooling makes the loops have lower and higher densities at high and low temperatures, respectively. The respective underdensity and overdensity at high and low temperatures have been reported in the analyses of coronal data by Winebarger *et al.* (2003).

5. Conclusions

The matter presented in the foregoing sections leads us to arrive at the following conclusions:

- The stressing models (including nanoflare scenario) involving slow footpoint motions of loops appear to be in a better position to explain coronal heating.
- Frequency distribution of nanoflare energies is a power-law with exponent α which must be greater than 2. The value of this exponent is still ambiguous and controversial.
- The actual observed time variation of the X-ray and EUV fluxes in the solar atmosphere may not be an evidence of small-scale magnetic reconnections (nanoflares) but could well be due to dissipation of Alfvén waves.
- The signatures of coronal heating processes should be searched in the transition region and lower corona because these regions have relatively smaller time scales.
- Only observations are capable of deciding whether the corona is primarily heated by nanoflare-like events or by Alfvén waves. For example, the Extreme Ultraviolet Imaging Spectrometer on Solar-B Satellite (to be launched in 2006) is capable of deciding whether the nanoflare events are actually MHD shocks due to Alfvén waves or not?

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Evolving Coronal Holes and Interplanetary Erupting Stream Disturbances

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Abstract. Coronal holes and interplanetary disturbances are important aspects of the physics of the Sun and heliosphere. Interplanetary disturbances are identified as an increase in the density turbulence compared with the ambient solar wind. Erupting stream disturbances are transient large-scale structures of enhanced density turbulence in the interplanetary medium driven by the high-speed flows of low-density plasma trailing behind for several days. Here, an attempt has been made to investigate the solar cause of erupting stream disturbances, mapped by Hewish & Bravo (1986) from interplanetary scintillation (IPS) measurements made between August 1978 and August 1979 at 81.5 MHz. The position of the sources of 68 erupting stream disturbances on the solar disk has been compared with the locations of newborn coronal holes and/or the areas that have been coronal holes previously. It is found that the occurrence of erupting stream disturbances is linked to the emergence of new coronal holes at the eruption site on the solar disk.

A coronal hole is indicative of a radial magnetic field of a predominant magnetic polarity. The newborn coronal hole emerges on the Sun, owing to the changes in magnetic field configuration leading to the opening of closed magnetic structure into the corona. The fundamental activity for the onset of an erupting stream seems to be a transient opening of pre-existing closed magnetic structures into a new coronal hole, which can support high-speed flow trailing behind the compression zone of the erupting stream for several days.

Key words. Sun: coronal holes—interplanetary disturbances—interplanetary scintillation.

1. Introduction

Coronal holes play an important role in the occurrence of various kinds of solar events. It is well established that the geomagnetic activity is related to the development of coronal holes on the Sun. Coronal transients, most of which are associated with eruptive prominences, have a significantly higher rate of occurrence in the vicinity of coronal holes than at other locations on the Sun (Webb *et al.* 1978). Active regions that produce solar flares and associated type II radio bursts are found to exist in the neighbourhood of coronal holes, indicating that MHD – shock waves generating type II – burst

emission migrate along open magnetic field lines of the coronal holes (Shelke & Pande 1985). Several large scale soft X-ray blowouts are seen to erupt from or near the boundary of coronal holes (Bhatnagar 1996). There are some observations that show that transient coronal holes (sometimes now called ‘X-ray dimming’ or “coronal depletions”) frequently form subsequent to eruptive prominences, coronal mass ejections or both (Rust 1983; Watanabe *et al.* 1992; Hudson 1996; Sterling & Hudson 1997).

The interplanetary scintillation (IPS) measurements at 81.5 MHz detect the enhancement in the density turbulence at distances greater than or equal to 0.7 AU. The large-scale structures of enhanced density turbulence in interplanetary medium are termed interplanetary disturbances. Hewish & Bravo (1986) used IPS measurements at 81.5 MHz and coronal hole images to study the solar cause of interplanetary disturbances. The proximity of mid-latitude coronal holes near the sources of interplanetary disturbances on solar disc led them to conclude that perhaps some coronal hole activity was to be blamed for interplanetary disturbances and interplanetary shocks.

In IPS maps of large-scale structures of higher or lower than average plasma density in the solar wind at 0.6–1.5 AU, Hewish & Bravo (1986) observed two kinds of typical interplanetary disturbances during August 1978–September 1979, namely: (1) most stable Co-rotating Interaction Regions (CIRs), and (2) transient erupting streams. The erupting stream disturbance has a roughly spherical compressed shell of enhanced plasma density, which is driven by the high-speed flow of low-density plasma trailing behind for several days. Hewish’s team mapped 73 erupting stream disturbances, derived their origins by back projecting the disturbances onto the solar disk and plotted their sources on the synoptic maps using Carrington coordinates as circles of about 90° wide. The positional agreement between coronal holes and 90° circles led them to conclude that mid-latitude coronal holes are the sources of erupting stream disturbances.

Although coronal hole is a source of an excess kinetic energy flux in the solar wind (Steinitz & Eyni 1980), it is difficult to understand how a stable coronal hole can trigger a transient erupting stream, unless it is evolving. A coronal hole evolves, owing to the changes in its magnetic field configuration. It is plausible that changes in the field configuration of an evolving coronal hole can onset the erupting stream. In this paper, an attempt has been made to investigate a possible association between evolutionary changes in coronal holes and sources of erupting streams on the solar disk to probe into this possibility. Here, section 2 describes the observational material and the subtraction method used to delineate evolutionary changes in coronal holes. These evolutionary changes are categorized in different types in section 3 and their association with erupting streams is established in section 4. The magnetic configuration of coronal regions overlying the eruption sites of erupting streams is examined in section 5. Section 6 discusses and concludes the results of the present study.

2. Observational data

It is necessary to identify evolutionary changes in coronal holes over which sources of erupting stream disturbances overlie, so as to establish the association between them. As the coronal hole evolves, its size increases or decreases due to opening or closing of its magnetic fields respectively. Owing to the changes in magnetic field configuration,

some parts of the evolving coronal hole disappear and/or are formed, that can be segregated by the ‘Subtraction Method’. In this method, coronal hole images seen in synoptic maps of two consecutive Carrington rotations are superimposed on each other to eliminate the coronal hole areas that are static (i.e., unchanged), and thereby to delineate the parts of coronal holes that are formed and/or disappeared from one synoptic map to another. These delineated parts represent the evolutionary changes in coronal holes.

Using the subtraction method, the areas of coronal holes which get diminished and/or are formed have been delineated and new synoptic charts, hereafter called “Evolution or E-charts”, are made by plotting these delineated parts of coronal holes in Carrington coordinates (cf. Fig. 1). Thirteen such E-charts are made for KPNO He I 10830 Å coronal hole images seen in synoptic maps of Carrington rotations 1671–1684 (August 1978–August 1979). Note that the E-chart does not incorporate static areas of coronal holes, as they will be eliminated in the subtraction. Data on coronal holes are taken from Solar Geophysical Data.

In the present study, about 68 events of transient erupting stream disturbances observed by Hewish & Bravo (1986) are analyzed. The data on their sources on the

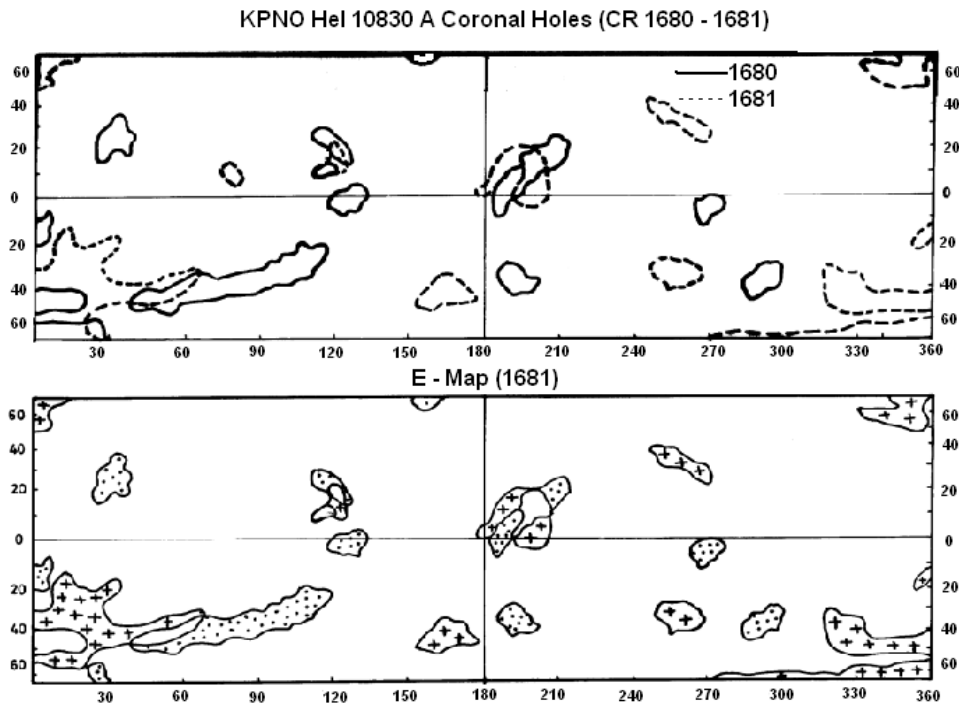


Figure 1. (Top) The superposition of coronal hole images seen in Carrington rotations 1680 and 1681. Solid and dashed lines represent the boundaries of coronal holes seen in rotations 1680 and 1681 respectively. (Bottom) The E-map (1681) derived from the above superposition. Crossed regions represent newly formed coronal hole areas (NFOCHA) formed due to the emergence of new coronal holes or the growth of existing coronal holes. Dotted regions correspond to the areas that have been coronal holes previously (POCHA). Note that E-chart does not incorporate static (i.e., unchanged) areas of coronal holes.

Sun are taken from Hewish & Bravo (1986). Other observed interplanetary disturbances are Co-rotating Interaction Regions (CIRs), which are forward–reverse MHD shock ensembles, together with the regions of enhanced density between them. They are produced in the interplanetary medium due to the interaction of fast solar wind streams with slower ones. Therefore it would seem that they do not originate on the Sun and hence they are not considered in this paper.

3. Evolutionary changes in coronal holes

The superimposition of coronal hole images seen in two consecutive rotations reveals that some coronal holes appear to grow bigger, and some appear to get smaller. Besides, the emergence of some new coronal holes and total disappearance of some existing coronal holes can also be seen. Interestingly, some existing coronal holes appear to grow to form new areas at the time when some areas of them are disappearing. These evolutionary changes in coronal holes have been classified into two categories, namely: (i) Previously Occupied Coronal Hole Area (POCHA) and (ii) Newly Formed Coronal Hole Area (NFOCHA). The former is related to the partial or complete disappearance of a coronal hole, whereas the latter is associated with the formation of new areas of coronal holes due to the birth of new holes or the growth of existing holes.

These two categories can be seen in Fig. 1. Here the regions filled with dots represent the areas of coronal holes, which existed in the preceding rotation 1680, but are seen to disappear in the succeeding rotation 1681. That is, these are the parts of coronal holes that have disappeared during the time between two synoptic maps. As they correspond to the locations that have been coronal holes previously, they are termed ‘Previously Occupied Coronal Hole Areas (POCHA)’. Similarly, regions filled with crosses are the areas of coronal holes, which did not exist in the synoptic chart of the preceding rotation 1680, but are seen to appear only in the synoptic chart of succeeding rotation 1681. In other words, they are the new areas of the coronal holes that are formed due to the emergence of entirely newborn coronal holes or the growth of coronal holes that existed in the preceding rotations. They are termed ‘Newly Formed Coronal Hole Area (NFOCHA)’.

4. Association of erupting stream disturbances with evolutionary changes in coronal holes

Hewish & Bravo (1986) have indicated the probable locations of the back projected erupting stream disturbances on the Sun as 90° wide circles on the synoptic maps. Using their data in the present study, the centers of these circles that ought to be the locations of origins of disturbances on the solar disk have been marked by bullets (●). These sources of erupting streams have been plotted on corresponding E-charts and their positions have been compared with the locations of delineated evolutionary changes in coronal holes in the form of POKHA and NFOCHA.

Distinct cases of newborn coronal holes over which the sources of erupting stream disturbances superimpose, have been identified. The classic example is erupting streams Nos. 7, 8 and 9 that came from the same source on rotation 1672, as indicated by the event numbers against the same Hewish’s circle (cf. Fig. 2d). The sections of 10830 Å-synoptic maps of rotations 1671 and 1672 reveal that an isolated new coronal

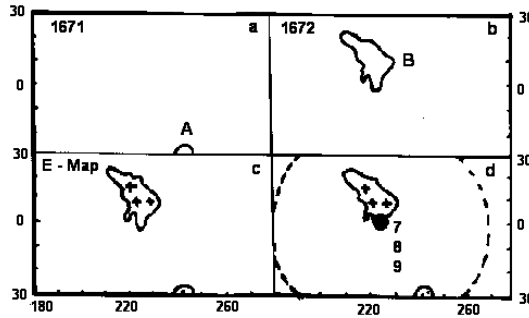


Figure 2. (a & b) Sections of 10830 Å synoptic maps of rotations 1671 and 1672, illustrating coronal holes A and B. (c) E-map (1672), derived from superposition of (a) and (b), shows a newborn coronal hole 'B' (crossed region) and a small area 'A' that has been coronal hole previously (dotted region). (d) The dashed-circle represents the back-projected erupting stream disturbance onto the Sun. The center of the circle, as shown by bullet (●), corresponds to the source of erupting stream on solar disc. The event numbers 7, 8 and 9 refer to Table 1 of Hewish & Bravo (1986). Note the existence of source of the disturbances on the newborn coronal hole 'B'.

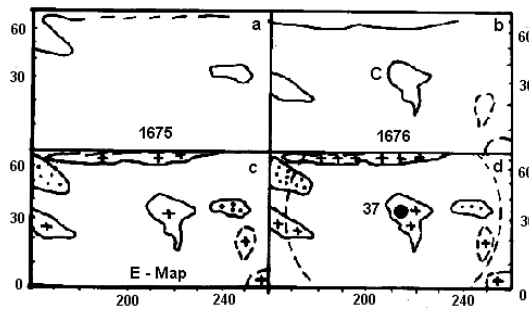


Figure 3. (c) E-map (1676) derived from superposition of (a) and (b) shows 5 newborn coronal holes (crossed regions) and 2 previously occupied coronal hole areas (dotted regions). (d) The source of back-projected erupting stream disturbance No. 37 on the solar disk, as shown by bullet (●), appears to exist on the newborn coronal hole 'C'. Note the existence of 4 NFOCHA and 2 POCHA within the circle that represents the back-projected disturbance onto the Sun.

hole 'B' (longitude 225–latitude N10; Fig. 2b) emerges on the solar disk in rotation 1672 and a small coronal hole 'A' (longitude 240–latitude S30) seen in rotation 1671 disappears subsequently in rotation 1672. The bullet (●) that indicates the source of the disturbances is seen to overlie on the newborn coronal hole 'B' (cf. Fig. 2d). It is interesting to note that there is no other coronal hole inside the Hewish's circle than the newborn coronal hole 'B' and a small area 'A' of about 10 degrees wide that has been a coronal hole previously (POCHA). However, 'A' is located far-off from the source location and thereby is unlikely to be linked to the erupting stream disturbances. Another example is erupting stream disturbance No. 37 (longitude 215–latitude N35; rotation 1676, Fig. 3d). The source of this disturbance appears to superimpose on the newborn coronal hole 'C' that emerges at longitude 230–latitude N35 in 1676 and resides with other 3 NFOCHA and 2 POCHA inside the Hewish's circle (cf. Fig. 3d). These examples demonstrate that the occurrence of erupting stream disturbances is related to the emergence of newborn coronal holes on the solar disk.

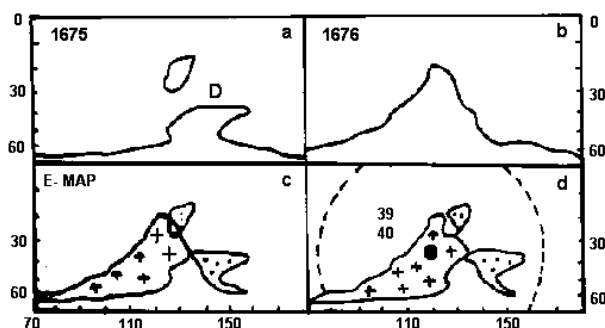


Figure 4. (a) and (b) Sections of 10830 Å synoptic maps showing polar coronal hole 'D' in preceding and succeeding rotations. (c) E-map (1676), showing concurrent growth and contraction of coronal hole D, forming NFOCHA (crossed region) on east side and POCHA (dotted region) on west side. (d) The source of back-projected erupting stream disturbances Nos. 39 and 40 on the Sun, as shown by bullet (●), appears to exist on newly formed coronal hole area. The dashed circle represents the back-projected disturbance on the Sun. Note the existence of both NFOCHA and POCHA within the circle.

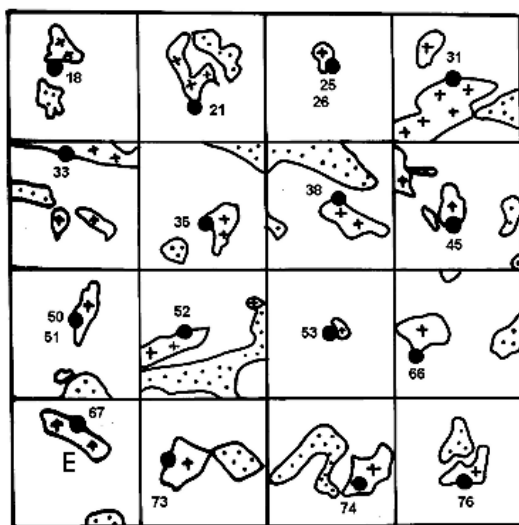


Figure 5. Striking examples of erupting stream disturbances demonstrating their association with NFOCHA (crossed regions). Dotted regions represent POCHA. Note the existence of sources of erupting streams (as shown by bullets) on NFOCHAs.

Additionally, erupting stream disturbances can also be related to new areas of coronal holes that are formed due to the growth of existing coronal holes. The best examples are erupting streams Nos. 39 and 40 that came from the same source located at longitude 120–latitude S35 in rotation 1676 (cf. Fig. 4d). Let us examine the evolution of a coronal hole 'D' (longitude 140–latitude S45; rotation 1675; Fig. 4a). It appears to grow eastward, but concurrently shrink westward in rotation 1676 (cf. E-map; Fig. 4c). Consequently, a new coronal hole area (crossed region) is developed into a large portion on the east side of 'D', when some of its west side area, shown by dotted region, is disappearing. The source location of erupting streams Nos. 39 and 40 appears to

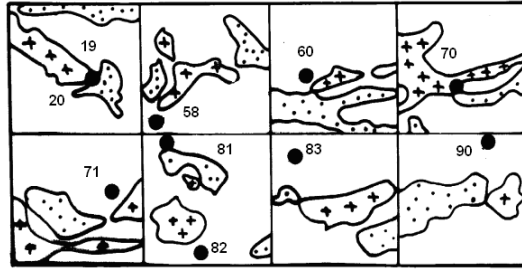


Figure 6. Classic examples of erupting stream disturbances showing a bit feeble association with NFOCHA. Note the existence of the sources of erupting streams (as shown by bullets) in close proximity to both NFOCHA (crossed regions) and POCHA (dotted regions).

exist inside this large, newly formed coronal hole area (NFOCHA) developed due to the growth of existing coronal hole (cf. Fig. 4d).

Some more striking examples are shown in Fig. 5. It can be seen that sources of erupting stream disturbances are located within the layout of NFOCHA. These NFOCHA cover both totally newborn coronal holes and new coronal hole areas formed due to the growth of existing coronal holes. All these examples demonstrate the association between source locations of erupting stream disturbances and newborn coronal holes on the solar disk (NFOCHA) formed due to the birth or growth of coronal holes.

However, there are some cases that exhibit a bit feeble association with NFOCHA. Figure 6 shows that the sources of erupting streams are located in close proximity to both NFOCHA and POCHA and thereby the possibility of their association with NFOCHA and POCHA is 50:50. The examples are event Nos. 19 and 20 (E-map 1673), 70 (E-map 1681) and 71 (E-map 1682). Besides, about 05 sources of the erupting stream disturbances appear to exist over or near the locations that have been coronal holes previously (POCHA) and hence are exceptions.

The histogram for the sources of erupting stream disturbances associated with NFOCHA or POCHA or both is plotted in Fig. 7. It can be seen that in a sample of 68 events of erupting streams, a large number of about 40 sources (plus 10 that are located within 20+ degrees as shown in Fig. 10) have shown their association with NFOCHA, with an exception of 5 events. About 13 sources have shown their alliance with both POCHA and NFOCHA. The sharp maximum for NFOCHA demonstrates a strong association between sources of erupting streams with newly formed coronal hole areas (NFOCHA). The critical examination of associated 40 NFOCHA reveals that all are not exclusively newborn coronal holes. Nearly half of them are new areas of coronal holes formed due to the growth of existing coronal holes. Their statistics are shown in Fig. 8.

There are some sources of erupting stream disturbances that do not overlies precisely on NFOCHAs, but reside close to them. This can be seen in Fig. 9. The classic example is event No. 64 (E-map 1680). The source is located about 10 degrees (in longitude) away from the border of the nearest NFOCHA. The histogram for the angular separation between the source locations and the border of associated NFOCHA or POCHA is shown in Fig. 10. Much closer positional agreement between sources of erupting streams and NFOCHA is apparent. This confirms the argument for a strong association between erupting stream disturbances and newborn coronal holes formed due to the birth or growth of coronal holes.

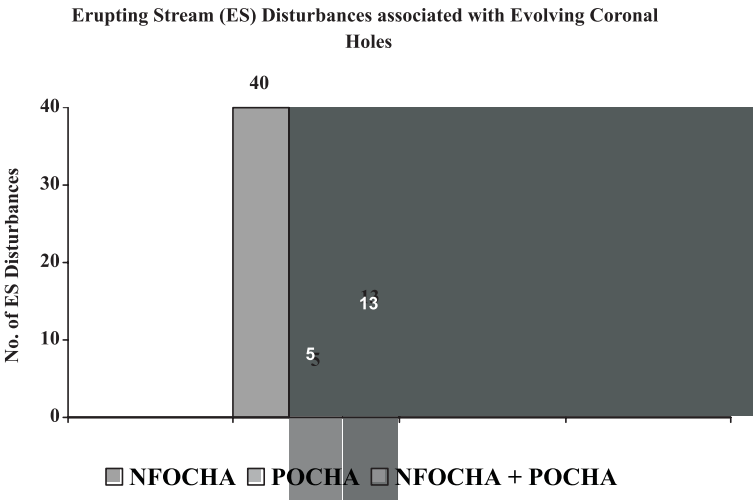


Figure 7. Histogram for sources of back-projected erupting stream disturbances associated with NFOCHA or POCHA or both. Note the sharp maximum for NFOCHA demonstrating their strong association with erupting stream disturbances.

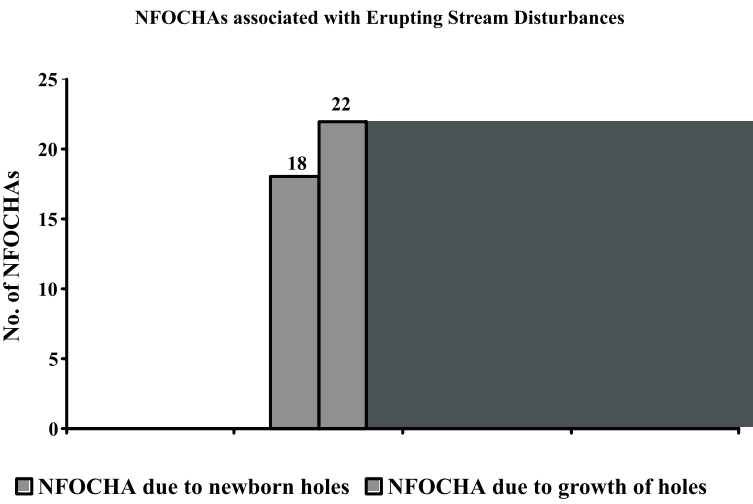


Figure 8. Statistics of newly formed coronal hole areas (NFOCHA) associated with erupting stream disturbances. About 18 NFOCHAs are exclusively newborn coronal holes and 22 are new areas of coronal holes formed due to the growth of existing coronal holes.

5. Magnetic configuration of coronal regions overlying the eruption site

It is interesting to examine the magnetic configuration of coronal regions in and around the eruption site of an erupting stream disturbance. This can be done by comparing daily maps of associated 10830 Å coronal holes before and after the date of eruption.

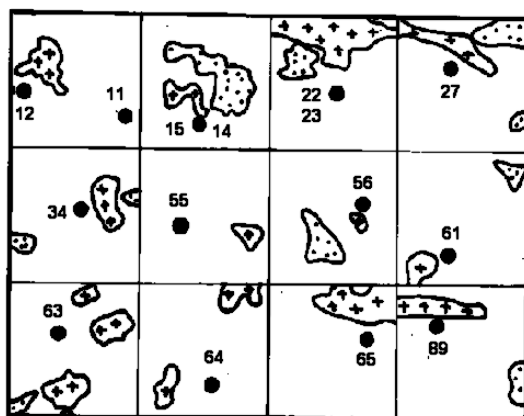


Figure 9. The classic examples of sources of back-projected erupting streams (shown by bullets) located close to NFOCHA (crossed regions) instead of overlying on them, as seen in preceding figures.

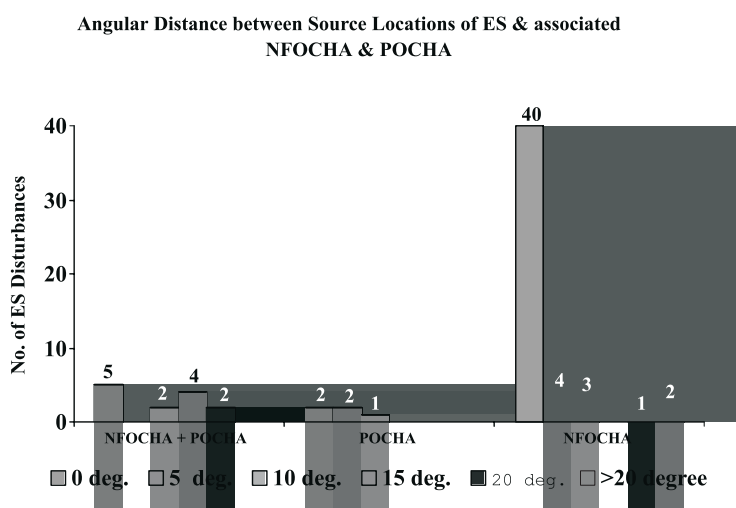


Figure 10. Histogram for sources of back-projected erupting stream disturbances (ES) associated with NFOCHA or POCHA or both for various angular separations (in degrees) between the source and border of nearest evolved part of coronal holes. Note the sharp maximum for NFOCHA for zero degree angular separation.

Let us examine an erupting stream disturbance No. 67 (cf. Fig. 5). This disturbance is unaccompanied by a flare or prominence activity anywhere on the disk. However, it is associated with the emergence of a new coronal hole 'E' on the Sun at latitude N 33–longitude 210 in rotation 1681. The daily maps of coronal hole 'E' arranged sequentially in Fig. 11 are examined to access the magnetic configuration of coronal regions in and around the eruption site of erupting stream disturbance No. 67. The critical examination of Fig. 11 reveals that coronal hole 'E' emerges on the solar disk at N 32 E 50

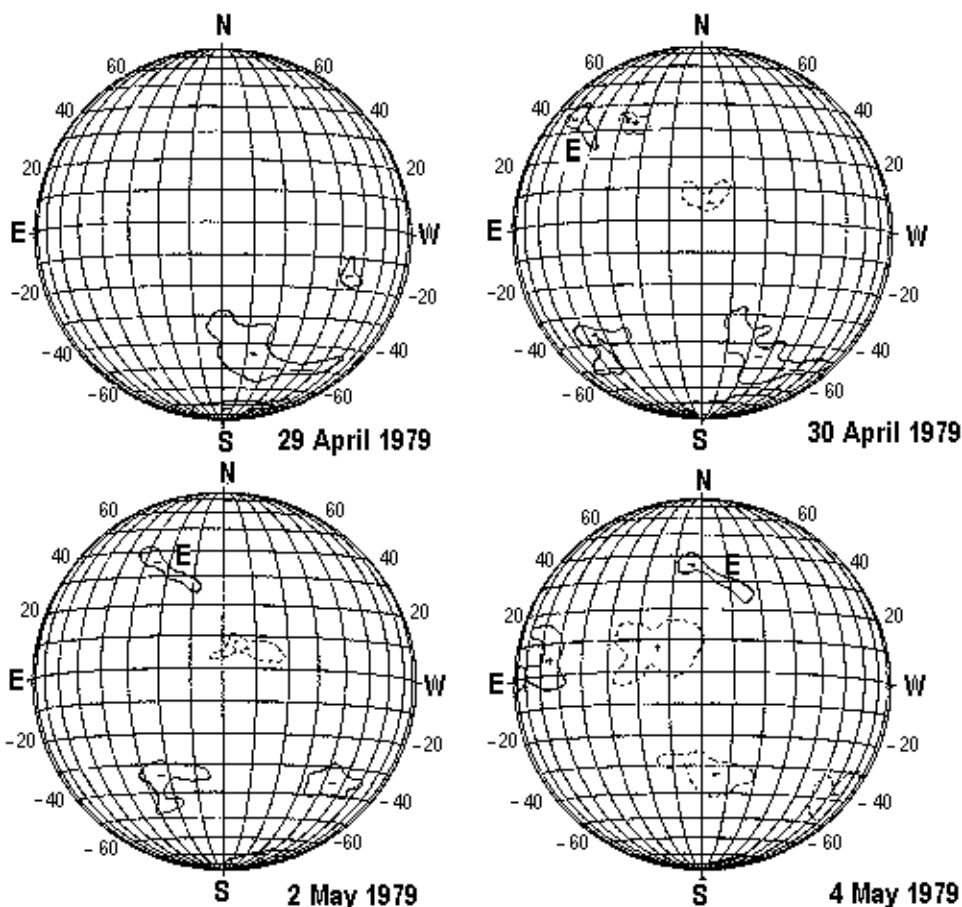


Figure 11. Day-to-day evolution of coronal hole 'E' associated with erupting stream No. 67. Note the formation of 'E' at N32–E50 on 30 April 1979. The arrival of leading edge of the erupting stream at 1 AU is 04 May 1979. There was neither flare nor disappearing prominence activity within the source area on the solar disc during a period $1 < \Delta t < 5$ days prior to the date of arrival.

on 30 April 1979 and can be seen in Fig. 11 till 4 May 1979. The leading edge of the disturbance has been reported to arrive at 1 AU on 4 May 1979 (cf. Hewish & Bravo 1986). The coronal hole 'E' should form within 90° -wide source area on the Sun during a period of 1–5 days prior to the arrival of the disturbance at 1 AU, if this newborn coronal hole was associated with erupting stream No. 67, which is indeed the case. This observation strengthens the argument for a close connection between transient erupting stream disturbances and newborn coronal holes lasting for days at the eruption site. It also suggests a conspicuous disruption of pre-existing closed magnetic structures in the corona during the onset of the erupting stream disturbance.

Like source location on the Sun, precise onset time of an erupting stream disturbance is not known. Therefore it is difficult to find out whether the evolution of coronal holes causes the eruption of erupting streams or eruption triggers the changes in the coronal holes. However, this observation categorically demonstrates that the formation

of new coronal holes and the onset of transient erupting streams are the manifestation of some mechanism that opens up the large-scale, closed magnetic structure overlying the eruption site.

6. Discussion and conclusions

Interplanetary disturbances are identified as an increase in the density turbulence compared with the ambient solar wind. Erupting stream disturbances are transient large-scale structures of enhanced density turbulence in the interplanetary medium driven by the high-speed flows of low-density plasma trailing behind for several days. The present study demonstrates a strong association between the emergence of newborn coronal holes on the solar disk and the occurrence of erupting stream disturbances in the interplanetary medium. The comparison of daily maps of associated coronal holes before and after the date of eruption of a disturbance confirms a conspicuous disruption of pre-existing closed magnetic structures overlying the eruption site in the corona during the onset of the disturbance.

A coronal hole signifies a region of radial magnetic fields. There must clearly be a predominant magnetic polarity for a coronal hole and perhaps a region of curl $B \neq 0$ at the photospheric boundaries of the hole. The association between newborn coronal holes and erupting streams is indicative of some mechanism that is responsible for both the phenomena. It is reasonable to conclude that the fundamental activity for the onset of an erupting stream seems to be a transient opening of closed magnetic structure of coronal fields into a newborn coronal hole, which can support high-speed flow trailing behind the compression zone of the erupting stream for several days. The spherical-shaped compression zone of an erupting stream is perhaps the 'curved-front' of the closed field structure that erupts into interplanetary medium during the opening process.

It is possible that the 90° -wide source circles may include active regions and strong solar eruptions, besides newborn coronal holes. For the erupting stream events studied in the present study, Hewish and Bravo (1986) found that

- (1) flares and disappearing prominences may, on some occasion, be related to erupting streams as peripheral events and
- (2) there are many occasions when erupting streams are unaccompanied by a flare or prominence activity anywhere on the disk.

For another data, Manoharan (1997) has shown that more than 80% of the interplanetary transients observed during 1986–1991 are associated with the eruptive prominences on the Sun, which are suggested to be the source of coronal mass ejections. Moreover, there are some observations that show that transient coronal holes frequently form subsequent to eruptive prominences, CMEs or both (Rust 1983; Watanabe *et al.* 1992; Hudson 1996; Sterling & Hudson 1997). In case of CMEs, the changes in the distribution of mass ejection latitudes with time do not correspond to those for solar features or activity related to small-scale magnetic structures such as sunspots, active regions, or H- α flares; they do resemble those of features related to large-scale magnetic structures, such as prominences and bright coronal regions (Hundhausen 1993). Thus, it is reasonable to conclude that erupting streams are generated either by several different mechanisms or alternately by some more fundamental process, which under suitable conditions, can also trigger eruptive prominences or CMEs and form new

coronal holes at the eruption site. The sequence of onset timings of erupting streams, eruptive prominences and newborn coronal holes needs to be examined by investigating these three types of events together.

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Mechanisms of Coronal Heating

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Abstract. The Sun is a mysterious star. The high temperature of the chromosphere and corona present one of the most puzzling problems of solar physics. Observations show that the solar coronal heating problem is highly complex with many different facts. It is likely that different heating mechanisms are at work in solar corona. Recent observations show that Magnetic Carpet is a potential candidate for solar coronal heating.

Key words. Magnetic field—corona—Magnetic Carpet—magnetic flux.

Researchers have been searching for the elusive heating of the corona ever since its million-degree temperature was measured more than half a century ago. Some invisible agent is spiriting out energy into the corona, pumping it up and keeping it distended. Magnetism most likely plays a pivotal role. An important realization from Michelson–Doppler Imager (MDI) on Solar and Heliospheric Observatory (SOHO) is that the quiet Sun is covered with small regions of mixed polarity, termed ‘Magnetic Carpet’ contributing to solar activity on a short time scale. Magnetic loops of all sizes rise far into the solar corona, arising from regions of opposite magnetic polarity in the photosphere. Energy released when oppositely directed magnetic fields meet in the corona is one likely cause for making the solar corona so hot.

The motion of magnetic flux concentrations, along with the continual appearance and disappearance of opposite polarity pairs of fluxes releases a substantial amount of energy. The process of emergence of an ephemeral region will drive reconnection with the overlying field, and the subsequent cancellation of collections of network elements in the network will drive large-scale reconnection between opposite-polarity elements. These processes have been shown convincingly to explain the appearance of X-ray bright points, which however, although important in their own right, make up only a small fraction of the total coronal heating (Priest *et al.* 2002). The second possible way may be the buffeting of the magnetic fragments by granulation on a time-scale of 10–50 s and at a velocity of 2–10 km s^{−1} which will generate magnetic waves that propagate up towards the corona and dissipate efficiently within the fibril coronal magnetic structure. The third reconnection will be driven in several ways in the corona, including braiding by the thrusting of footprints around one another (Parker 1979; Galsgaard & Nordlund 1996), binary reconnection between flux pairs (Priest & Schrijver 1999), and separate reconnection (Priest & Titov 1996; Longcope 1996, 1998) by the interaction of three or more sources at the separation curves that join

one null point to another. However, reconnection will also be driven at the myriads of separatix surfaces (Parker 1979, 1994) that separate from each other all the tiny narrow elementary coronal loops whose footprints are the intense flux tubes. The driving will be by the lateral relative motion, the fragmentation, and the cancellation of the minute magnetic fragments in the photosphere. This formation and dissipation of current sheets on separatix surfaces can be referred to as “flux-tube tectonics”. The heating is fairly uniform along the separatix so that each elementary coronal flux tube is heated uniformly.

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Spectroscopic Studies of Solar Corona

VI: Trend in Line-width Variation of Coronal Emission Lines with Height Independent of the Structure of Coronal Loops

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Abstract. We have obtained spectroscopic observations in coronal emission lines by choosing two lines simultaneously, one [Fe x] 6374 Å and the other [Fe xi] 7892 Å or [Fe xiii] 10747 Å or [Fe xiv] 5303 Å. We found that in 95 per cent of the coronal loops observed in 6374 Å, the FWHM of the emission line increases with height above the limb irrespective of the size, shape and orientation of the loop and that in case of 5303 Å line decreases with height in about 89 per cent of the coronal loops. The FWHM of 7892 Å and 10747 Å emission lines show intermediate behavior. The increase in the FWHM of 6374 Å line with height is the steepest among these four lines. We have also studied the intensity ratio and ratio of FWHM of these lines with respect to those of 6374 Å as a function height above the limb. We found that the intensity ratio of 7892 Å and 10747 Å lines with respect to 6374 Å line increases with height and that of 5303 Å to 6374 Å decreases with height above the limb. This implies that temperature in coronal loops will appear to increase with height in the intensity ratio plots of 7892 Å and 6374 Å; and 10747 Å and 6374 Å whereas it will appear to decrease with height in intensity ratio of 5303 Å to 6374 Å *versus* height plot. These findings are up to a height of about 200 arcsec above the limb. The varying ratios with height indicate that relatively hotter and colder plasma in coronal loops interact with each other. Therefore, the observed increase in FWHM with height above the limb of coronal emission lines associated with plasma at about 1 MK may not be due to increase in non-thermal motions caused by coronal waves but due to interaction with the relatively hotter plasma. These findings also do not support the existing coronal loop models, which predict an increase in temperature of the loop with height above the limb.

Key words. Sun: corona—coronal structures—emission line widths—techniques: spectroscopy.

1. Introduction

Temperature variations in the coronal loops may give a clue to the processes responsible to heat up the coronal plasma. The variation in FWHM of coronal emission lines occurs due to two reasons, one because of variation in temperature of the coronal plasma and the other due to variation in the non-thermal velocity. The intensity ratio of a suitable pair of coronal emission line yields information about the temperature of coronal plasma. The line-width of these coronal emission lines may give the non-thermal velocity of the plasma. Most of the time intensity ratios have been studied in EUV lines to determine the temperature and density variations in coronal loops using filter ratio technique (Winebarger *et al.* 2003). Earlier following the discovery (Tousey *et al.* 1973; Vaiana *et al.* 1973) that the hot plasma is contained in closed magnetic structures; Rosner *et al.* (1978) recognized that corona could be approximated as a set of one-dimensional loops that were thermally insulated from the surroundings. They solved the one-dimensional energy balance equation by assuming constant pressure, uniform heating, and a static, semi-circular loop with constant cross section, deriving two scaling laws that related the loop half-length and volumetric heating rate to the resulting high apex temperature and base pressure. Serio *et al.* (1981) extended the scaling laws to allow for non-uniform pressure. Porter & Klimchuk (1995) found that the line-of-sight emission measures of the loops observed with the Soft X-ray Telescope were typically not equal to the emission measures derived from the scaling laws. Kano & Tsuneta (1996) found that temperature and emission measure are the highest at the loop top, and decrease towards the foot-points. They also found that the correlation between the total energy loss and the gas pressure for the steady loops is consistent with the theoretical energy scaling law. From the intensity ratios of various coronal emission lines, Tsubaki (1975) showed the existence of temperature in-homogeneity in the coronal structures. In view of this fact we analysed the simultaneous spectroscopic observations made in coronal emission lines to study the variation of FWHM with height only (Singh *et al.* 1999, 2002, 2003a, 2003b). We found systematic increase in the FWHM of the 6374 Å line and decrease in the FWHM of 5303 Å emission line with height above the limb and small increase in the FWHM of 7892 Å and 10747 Å lines as compared to that of 6374 Å line. The systematic variations and surprising decrease in FWHM of the 5303 Å line with height motivated us to investigate the variation of the intensity ratio, also to determine the temperature variations with height above the limb and found complex variations in the line intensity ratio of coronal emission lines with height above the limb (Singh *et al.* 2004). In this paper we discuss the implications of our findings on the existing coronal loop models.

2. Observations and data analysis

Spectrographic observations of several coronal regions were obtained in the [Fe x] 6374 Å, [Fe xi] 7892 Å, [Fe xiii] 10747 Å and [Fe xiv] 5303 Å emission lines with 25-cm coronagraph at the Norikura Solar Observatory on several days during the years 1997–2004. The Coude type coronagraph provided us coronal image with image scale of 25'' per mm and 7 m focus Littrow type spectrograph formed spectrum with a dispersion of 2.17 Å per mm at the third order red. By choosing different orders for different coronal emission lines, we could observe 6374 Å line in the third order

and 5303 Å line in the fourth order simultaneously by using two CCD cameras. One CCD camera of 512×512 format with pixel size of 13.5×13.5 microns was directly mounted at the focal plan of the Littrow focus and was used to take spectrum around the red line. The second CCD camera of 1024×1024 format with pixel size of 24×24 microns coupled with a 20-cm aperture and 8.86 m focal length Cassegrain telescope was kept in the diffracted beam from the grating which over-spilled the Littrow mirror of the main spectrograph. This camera recorded the spectrum around the green line. Similarly, we could locate the Cassegrain telescope to obtain the red line spectra in the third order and both the infrared lines spectra in the second order simultaneously. We could also make observations in 6374 Å and 7892 Å emission lines simultaneously by locating the Cassegrain telescope at the proper place. The binned CCD camera provided us a dispersion of 31.8 m Å at the green line, 58.4 m Å at the red line, 25 m Å at the 7892 Å and 121 m Å per pixel at infrared lines. But the slit width of 160 microns mostly used to make observations restricted the spectral resolu-

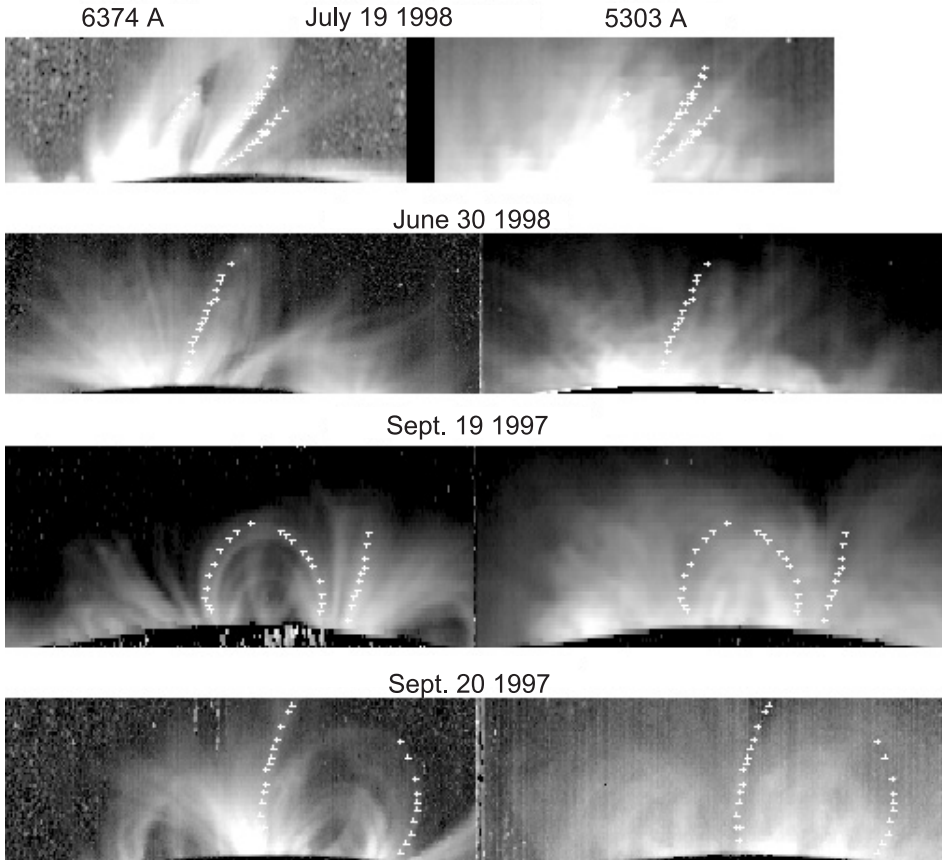


Figure 1. Intensity distribution in the coronal regions observed simultaneously in 6374 Å and 5303 Å coronal emission lines on different days. The + marks indicate the typical examples of different types of coronal loops. The variations in FWHM for both these lines with height above the limb are shown in Fig. 5 for these coronal loops.

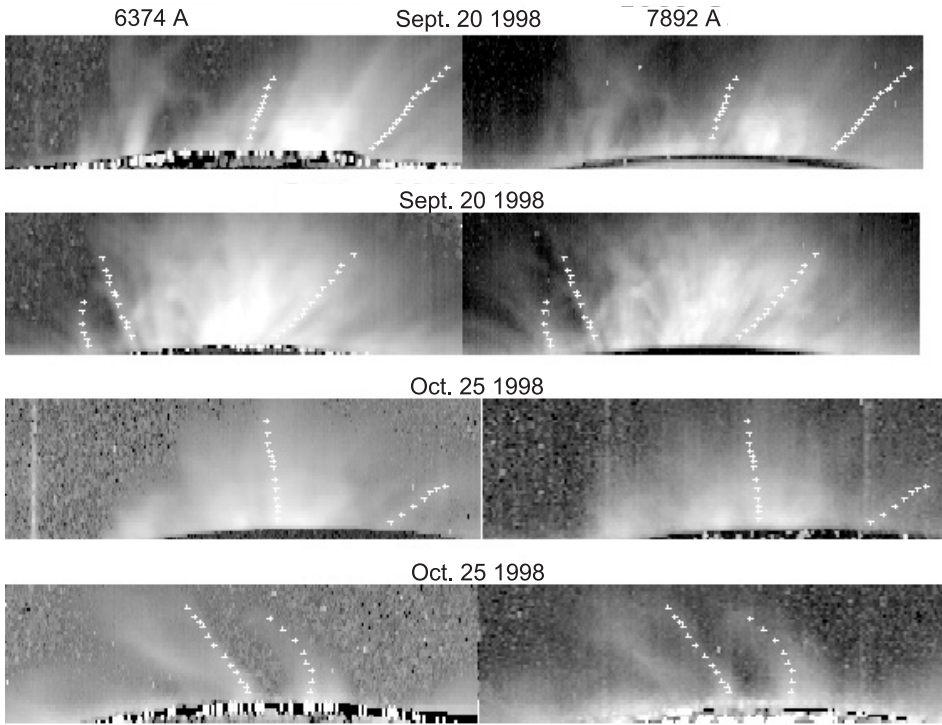


Figure 2. Same as for Fig. 1 but for 6374 Å and 7892 Å emission lines and variations in FWHM for the loops are shown in Fig. 6.

tions to 77 m Å, 128 m Å, 58 m Å and 291 m Å for the green, red, 7892 Å and infrared emission lines respectively and spatial resolution to 4". The typical exposure time was between 10 and 20 s for the red and green lines depending upon the sky conditions and between 20 and 70 s for the 7892 Å and infrared spectrum. The green and red line spectra were recorded with the same exposure times whereas exposure time for the red line was less compared to the 7892 Å and infrared lines. To maintain the simultaneity in observations, the exposures for both the CCD cameras were started at the same time and the CCD camera for recording the red line spectrum waited till the exposure for the infrared spectrum could be completed to begin the next exposure. The inclination of the glass block installed in front of the entrance slit was changed to obtain the successive spectra at different locations in the solar corona to build a two-dimensional image. One raster scan could be completed in 10–70 minutes depending on the exposure time and the number of steps. The typical raster scan consists of 50 positions with each step of 4" and covered $200'' \times 500''$ of the coronal region. All the spectra were corrected for the dark current, flat field and scattered light component due to sky brightness. A Gaussian fit was made to the observed emission line profile at each location of the observed coronal region to compute its peak intensity, central wavelength and width of the emission line. These values were then plotted as a function of spatial locations to obtain spectroheliograms, velocity-grams and width-grams in these coronal lines. The heliograms, which have different scales due to different

focal lengths of the optics used to image the spectra, were re-scaled and brought on the same spatial scale and dimensions using the solar disc spectrum with three fiducial marks. Figure 1 shows typical examples of heliograms of observed coronal regions on several days in the 6374 Å and 5303 Å lines simultaneously. Figure 2 shows the intensity distribution in some other coronal regions in the 6374 Å and 7892 Å and Fig. 3 shows the coronal regions observed in 6374 Å and 10747 Å emission lines simultaneously. To see the general trend in the parameters of emission lines, we selected 200–300 locations depending upon the area occupied by the coronal structures, with a spatial resolution of $4'' \times 4''$ on all the coronal structures visible in the red line intensity image. The corresponding locations in other images 5303 Å or 7892 Å or 10747 Å were selected automatically using software. Similarly we have done the analysis for individual coronal structures by selecting 10–20 locations depending upon its length to investigate the variations in emission line parameters as a function of height above the limb. To study the variation of line widths of these lines with height above the limb a linear fit to the FWHM data at various locations along the length of the individual coronal loop was made even though the linear fit may not be the best fit to the data.

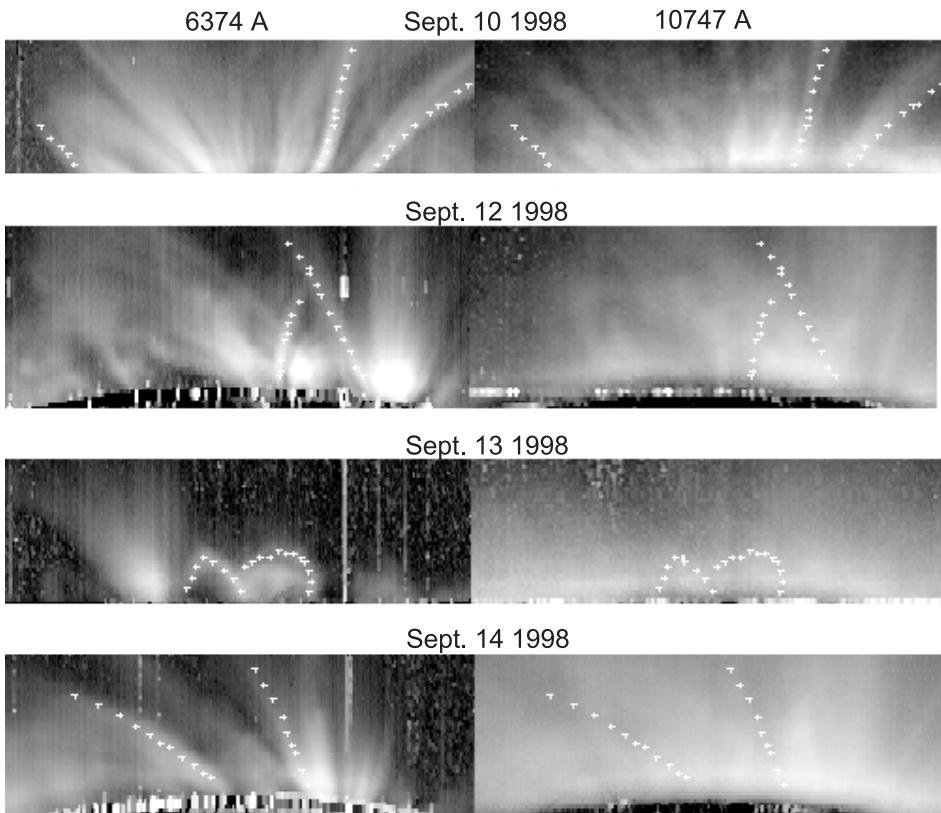


Figure 3. Same as for Fig. 1 but for 6374 Å and 10747 Å emission lines and variations in FWHM for the loops are shown in Fig. 7.

3. Results

Figure 4 shows that the line width of the [Fe XIV] 5303 Å emission line decreases with height above the limb in general and that of [Fe X] 6374 Å line increases with height above the limb. Line width of the [Fe XI] 7892 Å emission line also shows

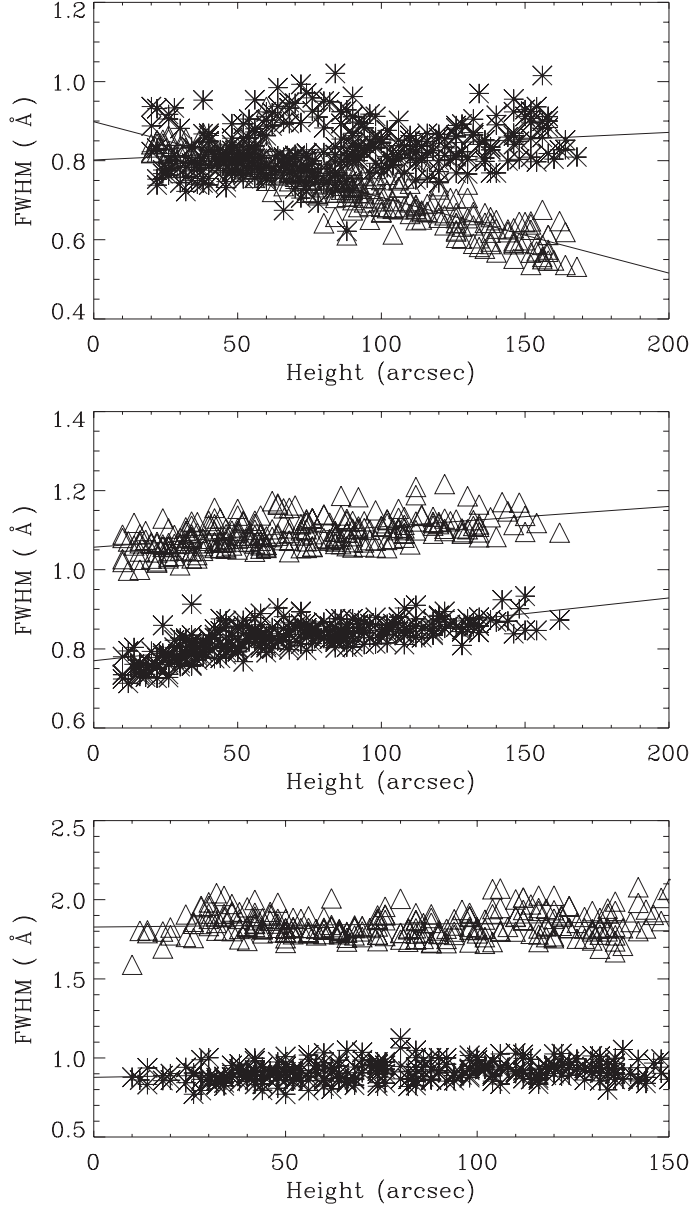


Figure 4. Top-most panel shows the variation of FWHM of the 6374 Å and 5303 Å emission lines with height above the limb when all the coronal loops in a coronal region are considered together. Middle panel shows the variations of 6374 Å and 7892 Å and bottom-most panel shows that of 6374 Å and 10747 Å emission lines.

increase with height but less steeply as compared to that for the 6374 Å line and the line width of [Fe XIII] 10747 Å shows marginal variations with height. We have analysed 1268 coronal loops observed in the 6374 Å line; 743 in the 5303 Å; 247 in the 7892 Å and 278 in the 10747 Å emission line. We found that 95% of the loops observed in 6374 Å line shows an increase in FWHM of the line with height above the limb; about 71% structures observed in 7892 Å line shows increase in FWHM with height; only 58% loops shows marginal increase in FWHM with height in 10747 Å line whereas 89% coronal loops shows decrease in FWHM with height above the limb in case of 5303 Å line. The increase in FWHM is maximum in case of 6374 Å emission line representing relatively low temperature plasma. Coronal loops of different sizes and shapes e.g., face-on and end-on; radial and non-radial; open and closed; were studied. Typical examples of the loops studied are shown by + marks in Figs. 1–3. Figure 5 shows the FWHM of the 6374 Å and 5303 Å emission lines

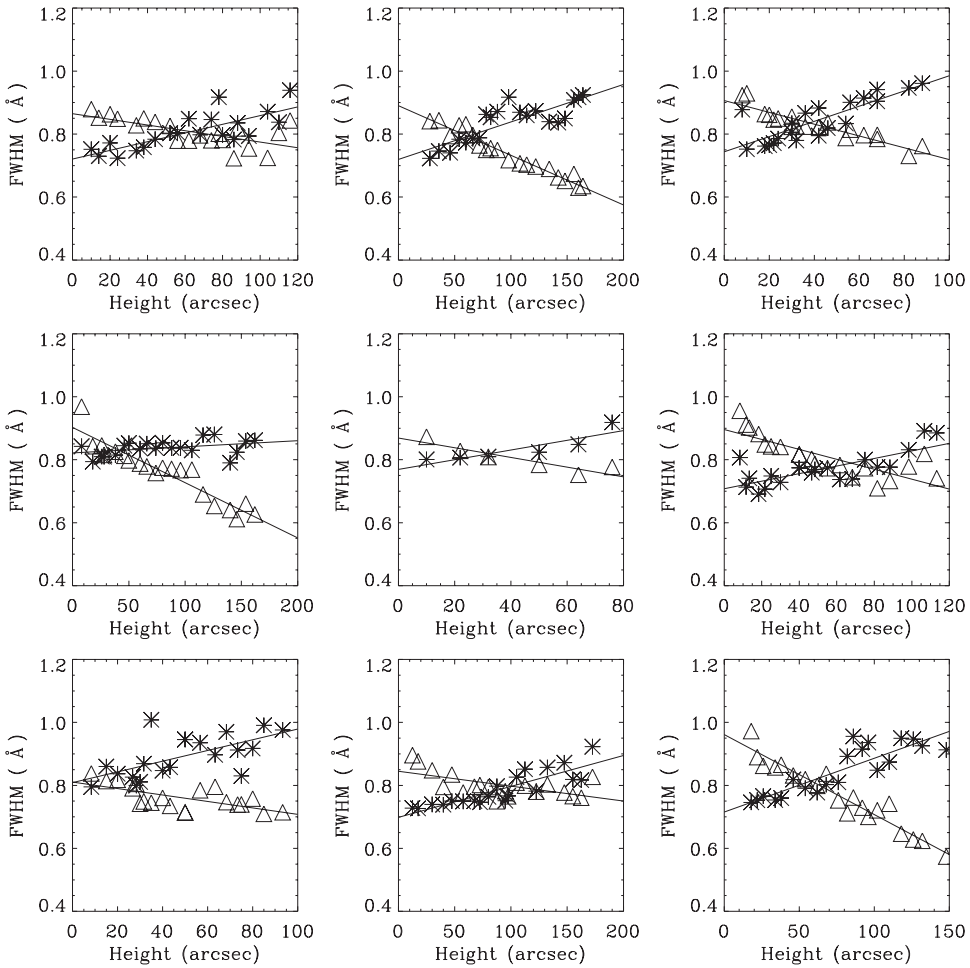


Figure 5. Nine panels show the variations of FWHM of 6374 Å and 5303 Å emission lines with height above the limb for various types of coronal loops shown in Fig. 1. The trend in variation of FWHM of 5303 Å line is the same in all the loops but the magnitude differs.

as a function of height in different types of loops. Figure 6 shows some other examples of the variation of FWHM in the 6374 Å and 7892 Å lines with height and Fig. 7 shows the same for 6374 Å and 10747 Å emission lines. The Figs. 5–7 indicate that the trend in the variations in FWHM with height for a chosen line is the same irrespective of the size, shape and orientation of the coronal loop. We also find that on an average, intensity ratio of 7892 Å to 6374 Å increases with height; that of 10747 Å to 6374 Å line remains almost unchanged with height and that of 5303 Å to 6374 Å decreases with height above the limb. Generally from the coronal loop models and abundances of ions as a function of temperature, one expects to observe a steeper increase in the intensity ratio of 5303 Å to 6374 Å as compared to that of 7892 Å to 6374 Å with height, opposite to the observed decrease in the intensity ratio with height.

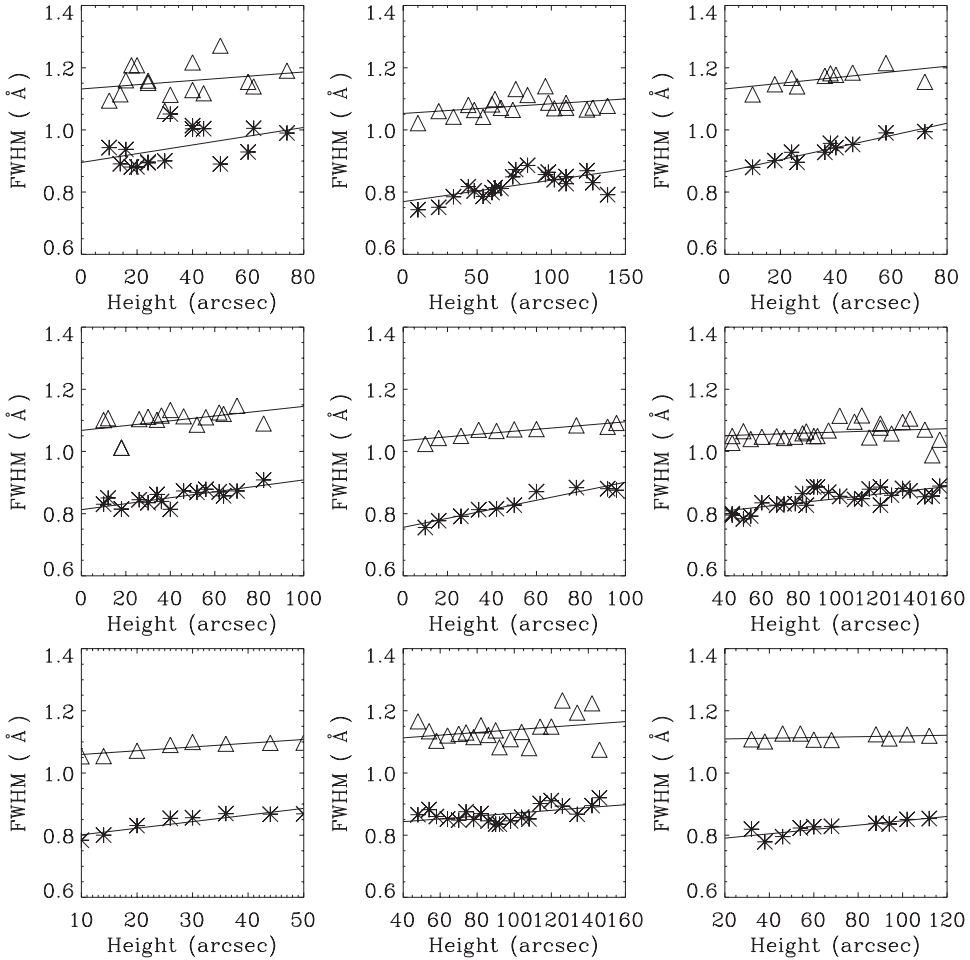


Figure 6. Same as that for Fig. 5 but for 6374 Å and 7892 Å emission lines and for coronal loops shown in Fig. 2.

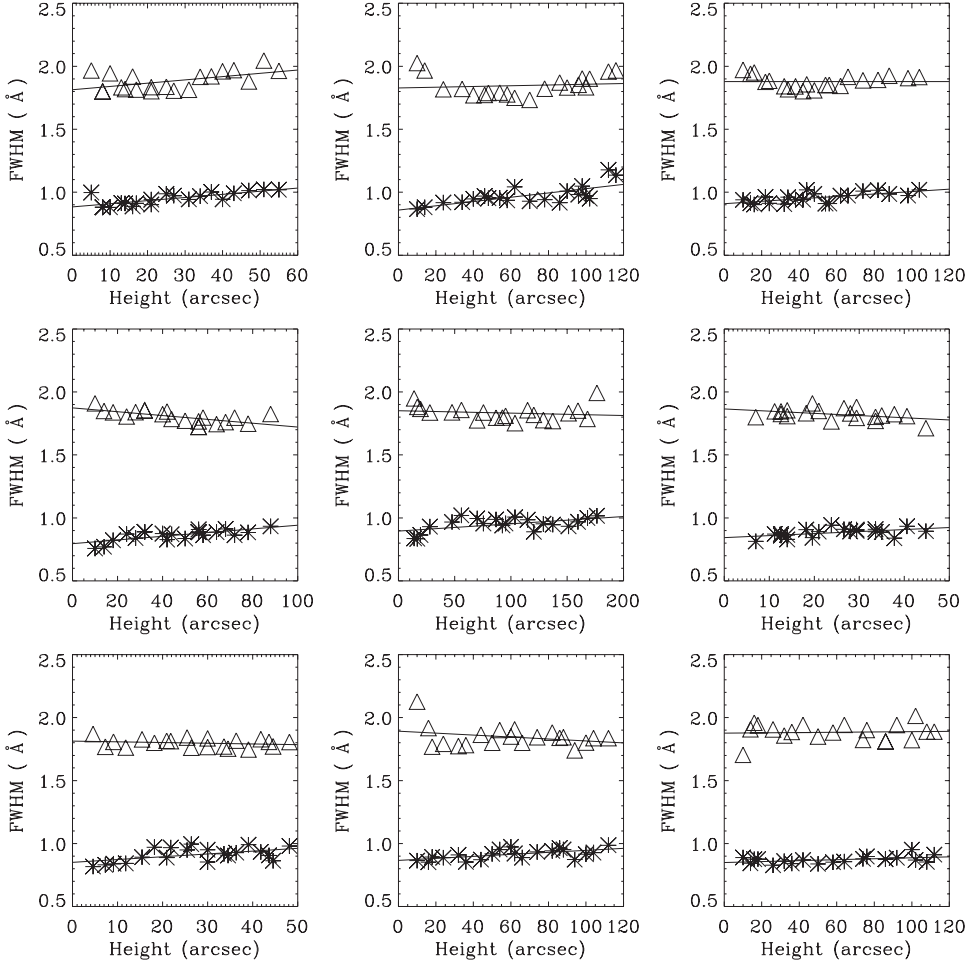


Figure 7. Same as that for Fig. 5 but for 6374 Å and 10747 Å emission lines and for coronal loops shown in Fig. 3.

4. Discussion and conclusion

These observations can be explained by assuming that the broad coronal loops connecting the magnetic field at the photospheric level and the fine structures in the corona are due to temperature and density variations in the broad loops and dynamic in nature. Probably plasma is fed in to these fine loops from the photosphere through chromosphere and the transition region and gets heated by small scale reconnections during its upward movement. The footpoints of the fine structures may be in different physical conditions leading to different temperature and density in the fine coronal loops. At larger heights the plasma in the fine structures interacts probably by collision and causes the relatively hotter plasma to share its energy with the relatively colder plasma. The effect of interaction of plasma goes on increasing with height above the limb. In the process, hotter plasma gets little cooler and cooler plasma gets little hotter and all the plasma in the coronal loop reaches a common average temperature around a

height of about 250 arcsec; the height depending on the activity and configuration of the coronal region and underlying magnetic fields. Under this assumption, we expect to observe an increase in the FWHM of the 6374 Å line and decrease in FWHM of 5303 Å line with height. We also expect to observe small variations in FWHM of 7892 Å and 10747 Å emission lines as these lines represent intermediate temperature plasma and thus will have smaller variations in temperature as compared to the variation of temperature in case of 6374 Å and 5303 Å line plasma. We also expect to observe an increase in the intensity ratio of 7892 Å to 6374 Å line as the temperature of plasma associated with both of these lines will increase in the loop with height and decrease in the intensity ratio of 5303 Å to 6374 Å line because temperature of the plasma representing 5303 Å line decreases with height. This scenario agrees with our findings. Therefore, the observed increase in FWHM with height above the limb of coronal emission lines associated with plasma at about 1 MK may not be due to increase in non-thermal motions caused by coronal waves but due to interaction with the relatively hotter plasma. These findings also do not support the existing coronal loop models, which predict an increase in temperature of the loop with height above the limb and put major restrictions on the loop models. New models need to be developed to explain the observations and those may include interaction of relatively hot and cold plasma.

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Physics of the Sun's Hot Atmosphere

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Abstract. This article briefly overviews the physics of the Sun's hot atmosphere, using observations from recent solar spacecrafts: Yohkoh, SOHO, TRACE and RHESSI.

Key words. Solar atmosphere—X-ray and EUV emission—coronal heating—solar wind.

1. Introduction

The Sun has well-defined regions which can broadly be compared to the solid part of an Earth-like planet and its atmosphere (cf., Fig. 1). Journeying out from the Sun's core, an imaginary observer will first encounter temperatures of 15 MK, where energy is generated by nuclear reactions. Temperatures get progressively cooler *en route* to the photosphere, a mere 6000 K. But then something entirely unexpected happens: the temperature gradient reverses. The chromosphere's temperature steadily rises to 10,000 K, and going into the corona, the temperature jumps to 1 MK. Parts of the corona associated with sunspots get even hotter.

How does the Sun's radiation, leaving the 6000 K surface, could heat the outer atmosphere to much higher temperatures? This question became and remains one of the central questions in astrophysics (Dwivedi & Phillips 2003). For comprehensive reviews on topics concerning the Sun's interior till its exterior and the solar observing facilities, the reader is referred to 'Dynamic Sun' (Dwivedi 2003). The first clue to the million-degree hot corona emerged in the 1940s when B. Edlén identified that the 'green' coronal line, observed during total eclipses, was due to 13-times ionized iron. This led to the two most basic properties of the corona, namely, that it is rarefied gas and that its temperature is very high. The discovery of soft X-ray and ultraviolet emission in the post-World War II years provided further proof of the corona's high temperature.

2. Solar X-rays and ultraviolet emission

Spacecrafts, built by the U.S. and Soviet space agencies in the 1960s and 1970s, dedicated to solar observations added much to our knowledge of the Sun's atmosphere; notably the manned NASA Skylab mission of 1973–1974. Ultraviolet and X-ray telescopes onboard gave the first high-resolution images of the chromosphere and corona and the intermediate transition region. Images of active regions revealed a complex of

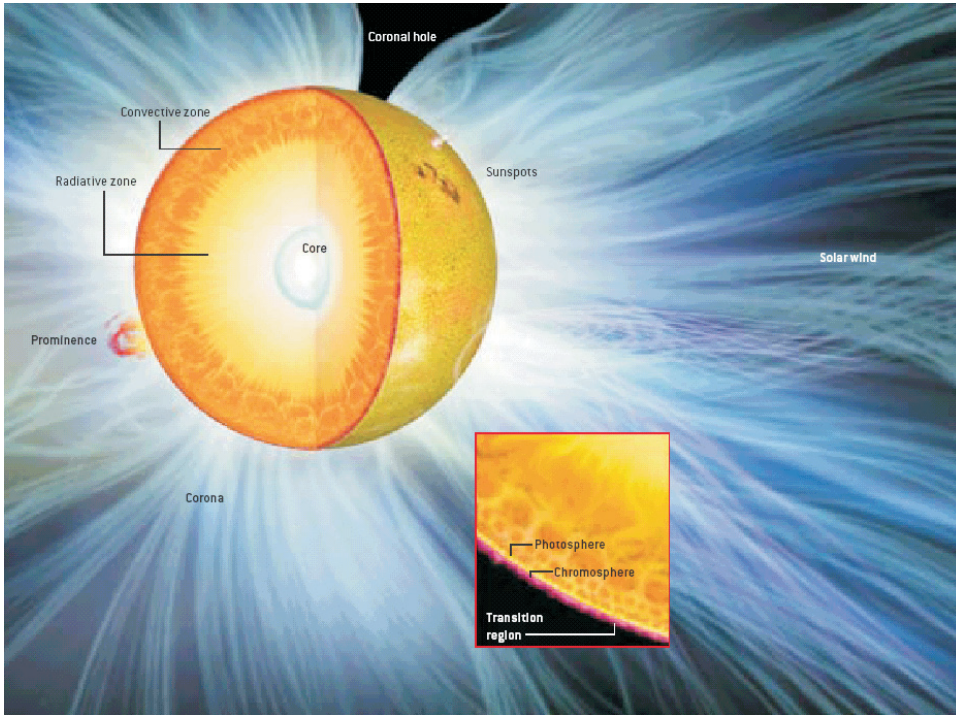


Figure 1. This figure gives a basic overview of the Sun's structure. The three major interior zones are the core (the innermost part of the Sun where energy is generated by nuclear reactions), the radiative zone (where energy travels outward by radiation through about 70% of the Sun), and the convection zone (in which convection circulates the Sun's energy to the surface). (Credit: Don Dixon.)

loops which varied greatly through their lifetimes, while ultraviolet images of the quiet Sun showed that the transition region and chromosphere followed the 'network' character previously known from Ca II K-line images. More recently, the spatial resolution of spacecraft instruments has steadily improved, nearly equal to what can be achieved with the ground-based solar telescopes (Mohan & Dwivedi 2001). Each major solar spacecraft since Skylab has offered a distinct improvement in resolution.

The ESA/NASA Solar and Heliospheric Observatory (SOHO), launched on December 2, 1995 into an orbit about the inner Lagrangian (L1) point, gives an uninterrupted view of the Sun (Dwivedi & Mohan 1997). There are several imaging instruments, sensitive from visible-light wavelengths to the extreme-ultraviolet. The Extreme-ultraviolet Imaging Telescope (EIT), for instance, uses normal incidence optics to get full-Sun images several times a day in the wavelengths of lines emitted by the coronal ions Fe IX/Fe X, Fe XII, Fe XV (emitted in the temperature range 600,000 K to 25,00,000 K) as well as the chromospheric He II 30.4 nm line. The Coronal Diagnostic Spectrometer (CDS) and the Solar Ultraviolet Measurements of Emitted Radiation (SUMER) are two spectrometers operating in the extreme-ultraviolet region, capable of obtaining temperatures, densities and other information from spectral line ratios. The Ultraviolet Coronagraph Spectrometer (UVCS) has been making spectroscopic observations of the extended corona from 1.25 to 10 solar radii from the Sun's centre,

determining empirical values for densities, velocity distributions and outflow velocities of hydrogen, electrons, and several minor ions. An intriguing observation with the UVCS has shown that particular ions, specifically highly ionized oxygen atoms, have temperatures in coronal holes, some 100 MK, that are much higher than those characterized by electrons and protons making up the bulk of the plasma (Kohl *et al.* 1997). There seems to be some directionality to the oxygen temperatures – they are higher perpendicular to the magnetic field lines than parallel to them, a result that is in agreement with interplanetary spacecraft sampling the particles making up the solar wind. The observation seems to call for high-frequency ion-cyclotron waves (Tu *et al.* 1998) that are produced lower down in the atmosphere as low-frequency waves but which through some cascade process end up at the observed frequencies. The excess broadening of coronal lines above the limb provides information on wave propagation in the solar wind. A striking difference in the width of line profiles observed on disk and in a polar coronal hole led to the discovery of the large velocity anisotropy, and consequently the possibility of the solar wind acceleration by ion-cyclotron resonance (Kohl *et al.* 1998).

The Transition Region and Coronal Explorer (TRACE) satellite went into a polar orbit around the Earth in 1998 (Handy *et al.* 1998). The spatial resolution is of order 1 arcsec (725 km), and there are wavelength bands covering the F IX, Fe XII, and Fe XV lines as well as the Ly-alpha line at 121.6 nm. Its ultraviolet telescope has obtained images containing a tremendous amount of small and varying features, for instance, active region loops are revealed to be only a few hundred kilometers wide, almost thread-like compared to their huge lengths. Their constant flickering and jouncing hint at the corona's heating mechanism. There is a clear relation of these loops and the larger arches of the general corona to the magnetic field measured in the photospheric layer. The crucial role of this magnetic field has been realized only during the past decade. The fields dictate the transport of energy between the surface of the Sun and the corona. The loops, arches and holes appear to trace out the Sun's magnetic field. The latest in the fleet of spacecraft dedicated to viewing the Sun is the Reuven Ramaty High Energy Solar Spectroscopic Imager (RHESSI), launched in 2002, which provides images and spectra in hard X-rays (wavelengths less than about 4 nm). RHESSI's observations of tiny microflares may provide clues to the coronal heating mechanism (Lin *et al.* 2002).

We can directly measure physical parameters such as electron density, temperature, flow speeds, etc. in the corona from emission line diagnostics. However, we cannot directly measure magnetic field strength, resistivity, viscosity, turbulence, waves, etc. New powerful tools of coronal seismology have enabled the detection of MHD waves by TRACE and EIT, spectroscopic measurements of line-widths by SUMER and CDS, ion and electron temperature anisotropy measurements with UVCS, and now microflares by RHESSI.

3. Coronal heating

The corona is a magnetically dominated environment consisting of a variety of plasma structures including X-ray bright points, coronal holes and coronal loops or arches (cf., Fig. 2). There is strong but indirect evidence to suggest that corona is heated by magnetic fields. One vital piece of information that we are still unable to measure is the corona's magnetic field strength. We can measure with considerable accuracy,

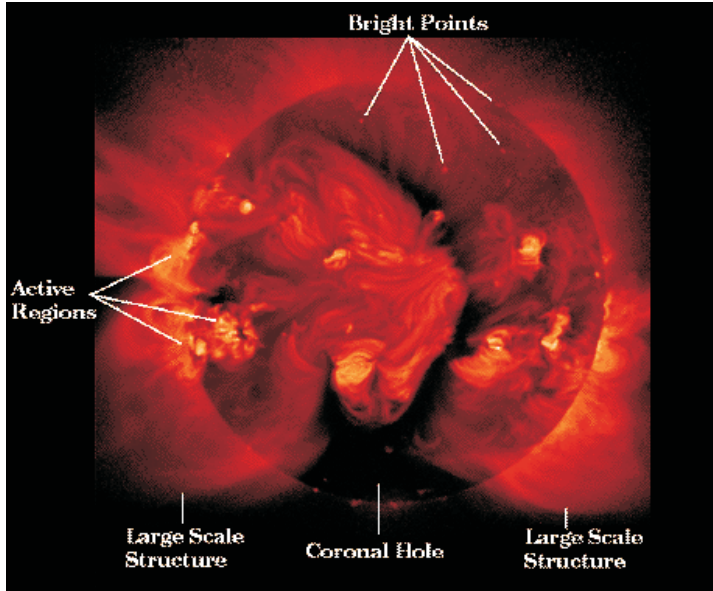


Figure 2. The X-ray Sun showing variety of plasma structures. (Credit: SXT/Yohkoh.)

the photospheric magnetic field using magnetographs that work on the Zeeman effect. Although, eventually, infrared measurements may give important information, practically the only way at present in which the coronal field can be deduced is through particular assumption: either that there are no currents flowing in the plasma (current-free), i.e., $\nabla \times \mathbf{B} = 0$, or that currents, if they exist, run parallel to magnetic field lines (force-free), i.e., $\mathbf{J} \times \mathbf{B} = 0$. These predict that the magnetic field of the corona generally has a strength of about 10 gauss. In active regions, the field may reach 100 gauss. It is to be noted, however, that vibrating magnetic structures in the solar atmosphere are opening up a new way to study the Sun. By fitting the observed oscillations from the TRACE spacecraft with MHD wave theory, Nakariakov & Ofman (2001) have determined the magnetic field strength in a coronal loop to be of about 13 gauss. The photospheric field in active region corona is more complex than in quiet regions and also the active region corona is appreciably hotter than in quiet regions. Thus there seems to be a relation between magnetic field strength, its complexity and heating.

A considerable problem with magnetic field heating is the fact that it requires the diffusion, and therefore reconnection, of magnetic field which implies a resistive plasma. However, the coronal plasma is, on the contrary, highly conducting. Using the induction equation of magnetohydrodynamics (MHD), we find that the diffusion time for a magnetic field is extremely long. If the characteristic distance over which diffusion occurs is as short as a few meters, then the diffusion time is a few seconds. In other words, the magnetic Reynolds number R_m , which measures how tied the magnetic field is to the plasma, is typically 10^{12} for the corona, indicating that the field is completely 'frozen in' to the plasma. However, reconnection requires R_m to be very small, much less than one. Thus, only if the length scales are very small (a few meters), can one achieve magnetic reconnection. Very small length scales do occur in the region of neutral points or current sheets, where there are steep magnetic field gradients which give

rise to large currents. It is thought, then, that such geometries are important for coronal heating if this is by very small energy releases, known as nanoflares. Some 10^{16} J are released in a nanoflare, which is 10^{-9} of a large solar flare. Many energy releases like this occurring all over the Sun, quiet as well as active regions, could account for heating of the corona. However, it is doubtful whether this mechanism would apply to coronal hole regions where the field lines are open to interplanetary space (for more details, see Phillips & Dwivedi 2003).

The above reasoning applies equally to the competing wave heating hypothesis of coronal heating, in which MHD waves generated by photospheric motions (e.g., granular or supergranular convection motions) are damped in the corona. In this case, we need conditions such that the magnetic field changes occur in a shorter time than, say the Alfvén wave transit time across a closed structure like an active region or quiet Sun loop. The waves, generated by turbulent motions in the convection zone or at the photosphere, may be surface waves, or body waves which are guided along the loops and are trapped. Theoretical work shows that short-period fast-mode and slow-mode waves (periods less than 10 s) could be responsible for heating (Porter *et al.* 1994) since the damping rates are high enough only for them. It is possible that (and would certainly be simpler) if X-ray bright points, coronal holes and the different types of coronal loops (see Fig. 2) are heated by the same mechanism, but it is also possible that they are heated by different mechanisms. The debate on coronal heating centres on whether the field energy is dissipated via numerous small magnetic reconnections ('nanoflares'), or damping of MHD waves which emerge from the interior and pass through the Sun's surface layers to the corona.

The over-riding problem for both theories is the way reconnections or MHD waves manage to dissipate current energy into heat energy in view of the fact that the corona is highly conducting. TRACE observations and analysis of damped coronal magnetic loop oscillations provide evidence that the resistivity of the corona is not in fact as low as it would seem classically (Nakariakov *et al.* 1999). But rather, there are anomalous processes in which heat dissipation occurs via particle-wave interactions instead of particle-particle collisions. If the resistivity is enhanced (or the resistivity and the viscosity are enhanced), then it would support the dissipation of waves and magnetic reconnection mechanisms for coronal heating.

Although the weight of observational evidence generally points to a nanoflare mechanism for providing the bulk of coronal heating, MHD waves probably contribute significantly. It is, for example, unlikely that nanoflares could have much effect in coronal holes. In these regions, the field lines open out into space rather than loop back to the Sun, so a reconnection would accelerate plasma rather than heat it. Yet the corona in holes is still hot. Astronomers have therefore scanned coronal holes as well as closed field regions for signatures of wave motions, which may include periodic fluctuations in brightness. The difficulty is that the MHD waves involved in heating probably have very short periods, just a few seconds. At present, spacecraft imaging is too sluggish to capture them. For this reason ground-based instruments operating during total eclipses are still important.

4. Wave activity

Hassler *et al.* (1990) carried out rocket-borne experiments to observe off-limb linewidth profile of Mg X 609 and 625 Å and reported the increasing linewidth with altitude

(up to 70,000 km). This observation provided the signature of outward propagating undamped Alfvén waves. SUMER and CDS spectrometers on SOHO recorded several line profiles and found the broadening of emission lines. These results are consistent with outward propagation of undamped Alfvén waves travelling through regions of decreasing density (Erdélyi *et al.* 1998). Harrison *et al.* (2002) reported the narrowing of the Mg X 62.50 nm line with height in the quiet near-equatorial solar corona, and concluded that this narrowing is a likely evidence of dissipation of Alfvén waves in closed field-line regions. Similarly, a significant change in slope of the linewidth as a function of height was seen in polar coronal holes by O'Shea *et al.* (2003) at an altitude of ≈ 65 Mm. These results obtained with the CDS spectrometer, if confirmed, could be crucial in understanding the coronal heating mechanisms. Due to the broad instrumental profile, the CDS instrument can only study line-width variations and cannot provide measurements of the linewidth itself, and, hence, of the effective ion temperature. Since the latter quantity is critical in constraining theoretical models of coronal heating and solar-wind acceleration, for instance, through the dissipation of high-frequency waves generated by chromospheric reconnection, Wilhelm *et al.* (2004) studied the problem further by analysing data recorded with the SUMER spectrograph in the Mg X doublet together with other neighbouring lines in both the quiet equatorial corona and in a polar coronal hole. Due to the high spectral resolution of SUMER, they were able to obtain profiles of both Mg X emission lines and measure their widths and variations as a function of height. Their work showed that linewidths of both components of the Mg X doublet measured by SUMER monotonically increase in the low corona in equatorial regions in altitude ranges for which scattered radiation from the disk does not play a major role. They did not find any evidence for a narrowing of the emission lines above 50 Mm. The same statement applies for a coronal hole, but they could not exclude the possibility of a constant width above 80 Mm (cf., Fig. 3).

While there have been reports of emission-line broadening with altitude, using SUMER, the CDS observations presented by Harrison *et al.* (2002) appeared to show emission-line narrowing. In order to resolve the apparent discrepancies, a joint CDS/SUMER observational sequence was successfully executed during the SOHO/MEDOC campaign in November and December 2003. The joint measurements were performed near the east limb (cf., Wilhelm *et al.* 2005). The pointing locations of the spectrometers are shown in Figs. 4(a) and 4(b) superimposed on He II and Fe XII solar images taken by EIT (Wilhelm *et al.* 2005). In the western corona, SUMER made additional exposures with a similar observational sequence from 13:02 to 15:06 UTC. Since the corona was very hot there, the slit positions are shown together with the Fe XII and Fe XV windows of EIT in Fig. 4(c) and 4(d).

The relative widths of the spectral lines observed by SUMER are shown in Fig. 5, separately for the eastern and western FOVs. In cases for which more than one observation was available, mean values have been shown. The range of $v_{1/e}$ extends from 35 km s^{-1} to 49 km s^{-1} in the relatively quiet eastern corona, and from 33 km s^{-1} to 48 km s^{-1} in the active western corona. No significant differences could be noticed within the uncertainty margins. The Mg X and Ca X lines show slight increases with height in the east, but are rather constant in the west. The Fe XII line is a little narrower in the west. Of particular interest was that the lines of Ca XIII and Fe XVIII were seen in the west, where they could be compared with the Ca X and Fe XII lines along the same LOS. In the relatively quiet equatorial corona above a small prominence, Wilhelm *et al.* (2005) found none or very slight increase of the linewidths of coronal emission lines

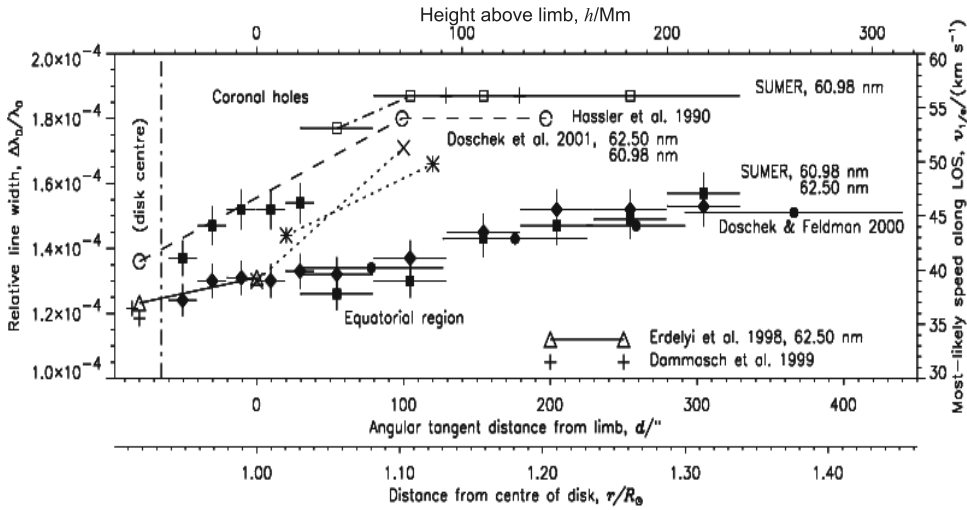


Figure 3. Relative linewidth variations as a function of radial distance from the Sun. Literature data and results obtained in this work (annotated ‘SUMER’) are compiled for the Mg X doublet in equatorial and coronal-hole regions. A classification of the Hassler *et al.* observations, of which only three typical values are shown, is not defined in their 1990 paper. An approximate scale of 715 km has been used for both the SOHO observations and those from the Earth. Integration intervals and the SUMER uncertainty margins are marked by horizontal and vertical bars. Related data points are in some cases connected by lines of various styles. These are meant to improve the orientation for the reader, but not as physical interpolations, in particular, for those points representative of centre-of-disk values displayed here near $-80''$ (from Wilhelm *et al.* 2004).

with altitude from measurements both with CDS and SUMER. Taking the combined uncertainty margins into account, the relative variations for Mg X were considered to be consistent, although the absolute widths could not be compared given the different instrumental spectral transfer functions. The SUMER observations indicated even less line-width variations with height in the more active corona. Singh *et al.* (2003) found height variations of line profiles in the visible light that depended on the formation temperatures of the lines. So, the solar conditions appear to have a direct influence on the line-width variations with height. Whether this dependence can account for the past apparent discrepancies cannot be decided unambiguously with the information available. More observations for different coronal activity levels are needed for this task. However, this joint study concluded that CDS and SUMER relative line-width measurements did not lead to inconsistencies if the same solar region was under study.

5. Coronal holes and the solar wind

Observations from the Skylab firmly established that the high-speed solar wind originates in coronal holes which are well-defined regions of strongly-reduced ultraviolet and X-ray emissions (Zirker 1977). More recent data from Ulysses show the importance of the polar coronal holes, particularly at times near the solar minimum, when dipole field dominates the magnetic field configuration of the Sun. The mechanism for accelerating the wind to the high values observed, of the order of 800 km s^{-1} , is not yet fully understood. The Parker model is based on a thermally-driven wind. To

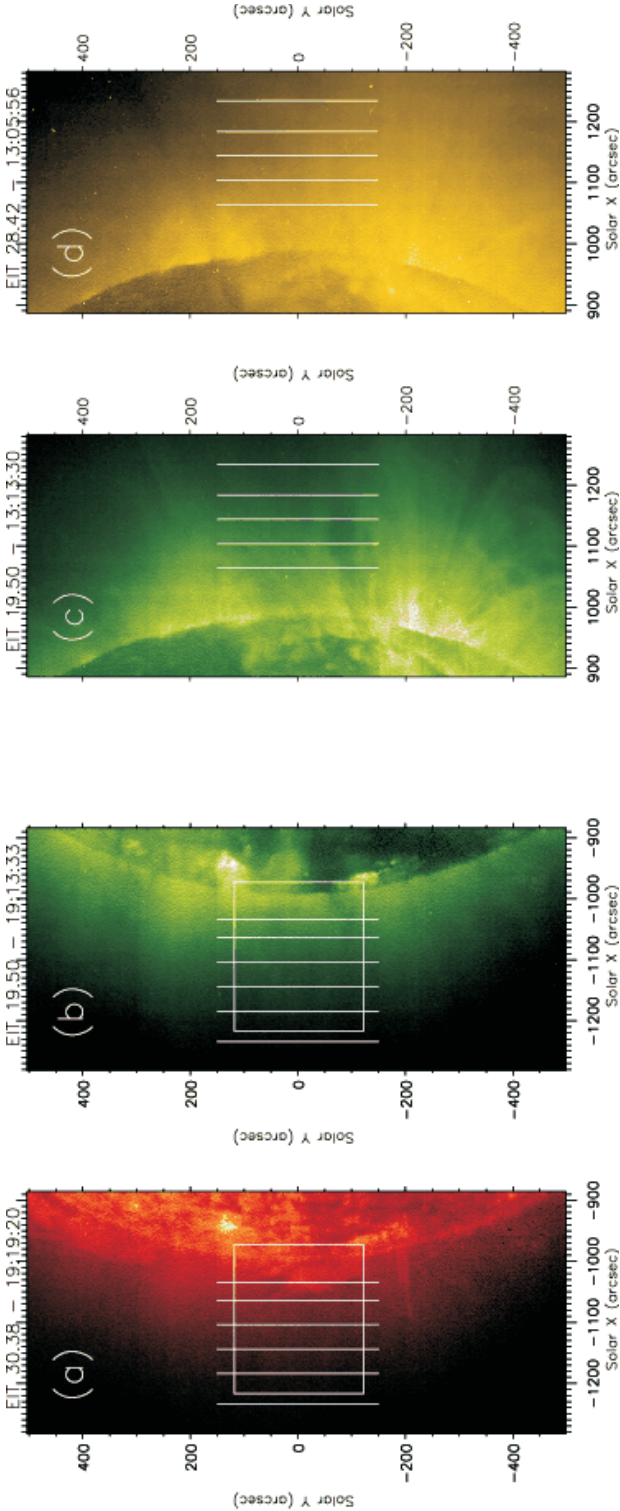


Figure 4. Positions of the CDS FOVs (rectangles) and the SUMER slit pointing locations in relation to He II, Fe XII and Fe XV solar images of 4 December 2003 (courtesy of the EIT consortium). Joint observations were obtained in the eastern corona. At low altitudes, a prominence caused a slight disturbance there. (a) He II spectral window; and (b) Fe XII window near the east limb; (c) Fe XII window; and (d) Fe XV window near the west limb (from Wilhelm *et al.* 2005).

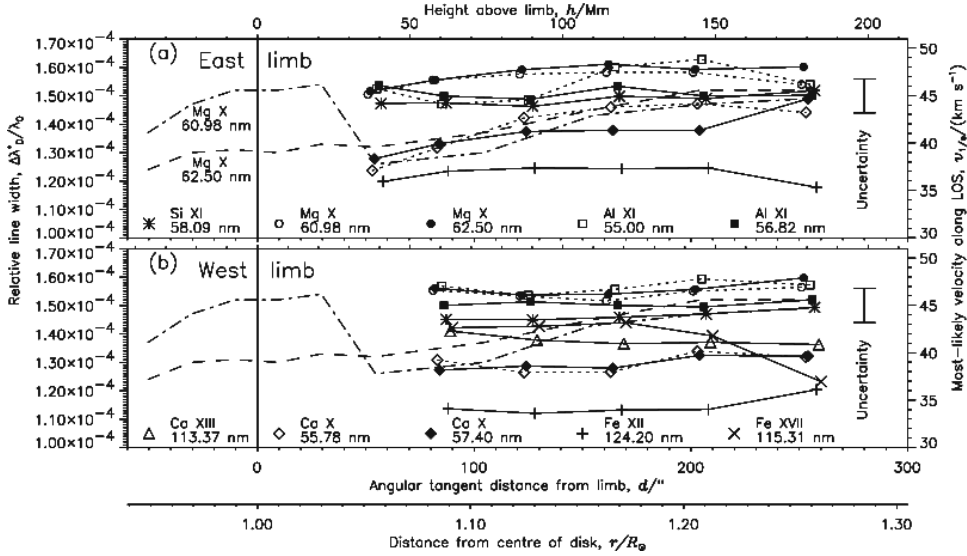


Figure 5. Summary of the SUMER linewidth measurements in terms of the relative linewidth, $\Delta\lambda_D^*/\lambda_D$, of the various ions, as well as their most-likely velocity along the LOS, $v_{1/e}$, as a function of distance from the limb. **(a)** Measurements above the east limb; **(b)** above the west limb. The dashed and dash-dotted lines indicate the Mg X observations of SUMER in November 1996. The increase of the width of the 60.98 nm line near the limb is caused by blends of O III and O IV transition-region lines (cf. Wilhelm *et al.* 2004). Near the centre of the solar disk a value of $\Delta\lambda_D^*/\lambda_D \approx 1.2 \times 10^{-4}$ was observed for Mg X (Erdélyi *et al.* 1998). At each slit position, an altitude range of $\approx 10''$ was covered. In the interest of clarity, we spread the plot symbols of the various spectral lines over this range (from Wilhelm *et al.* 2005).

reach such high velocities, temperatures of the order 3 to 4 MK would be required near the base of the corona. However, other processes are available for acceleration of the wind, for example, the direct transfer of momentum from MHD waves, with or without dissipation. This process results from the decrease of momentum of the waves as they enter less dense regions, coupled with the need to conserve momentum of the total system. If this transfer predominates, it may not be necessary to invoke very high coronal temperatures at the base of the corona.

Prior to the SOHO mission, there was very little information available on the density and temperature structure in coronal holes. Data from Skylab was limited, due to the very low intensities in holes and poor spectral resolution, leading to many line blends. High-resolution ultraviolet observations from instruments on SOHO spacecraft provided the opportunity to infer the density and temperature profile (see Fig. 6a) in coronal holes (Wilhelm *et al.* 1998). Comparing the electron temperatures with ion temperatures, it was concluded that ions are extremely hot and the electrons are relatively cool. Using the CDS and SUMER instruments on the SOHO spacecraft, electron temperatures were measured as a function of height above the limb in a coronal hole. Observations of two lines from the same ion, O VI 1032 Å from SUMER and O VI 173 Å from CDS, were made to determine temperature gradient in a coronal hole (David *et al.* 1998). This way temperature of around 0.8 MK close to the limb was deduced, rising to a maximum of less than 1 MK at $1.15 R_\odot$ (see Fig. 6b), then falling

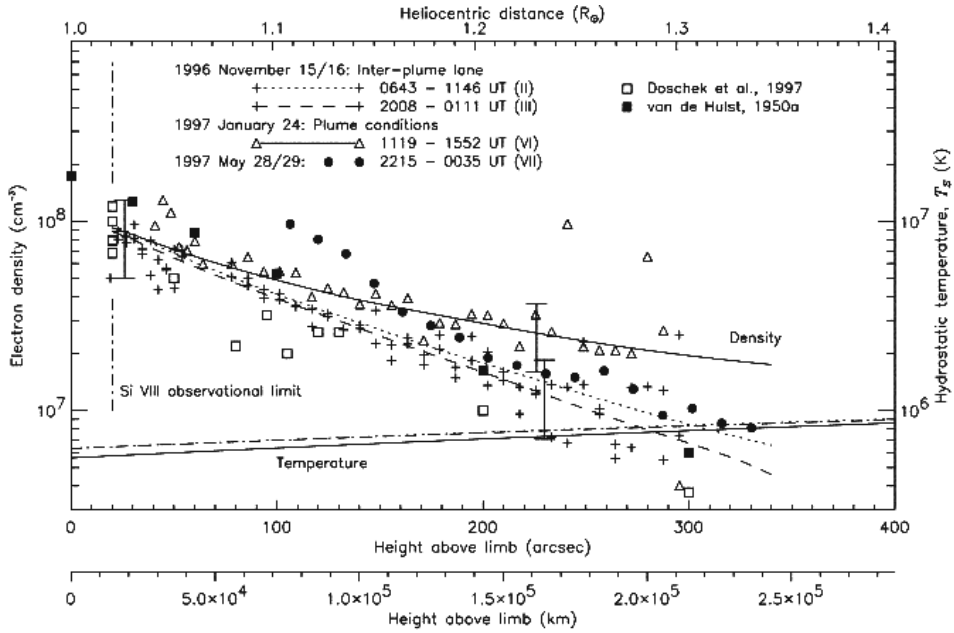


Figure 6(a). Electron densities derived from Si VIII line ratios and a comparison with data in the literature. The hydrostatic temperature, T_s , used for the fits of the line Si VIII ($1445 \text{ \AA}$) is plotted in the lower portion of the diagram, with a scale on the right-hand side. The points labeled '1997 May 28/29' are obtained from the ratios observed west of the polar plume assembly in a very dark region of the corona. The error bars indicate a density variation resulting from a $\pm 30\%$ uncertainty in the line ratio determination. Note that there is a small (3%) seasonal variation between the angular and the spatial scales for different data sets (from Wilhelm *et al.* 1998).

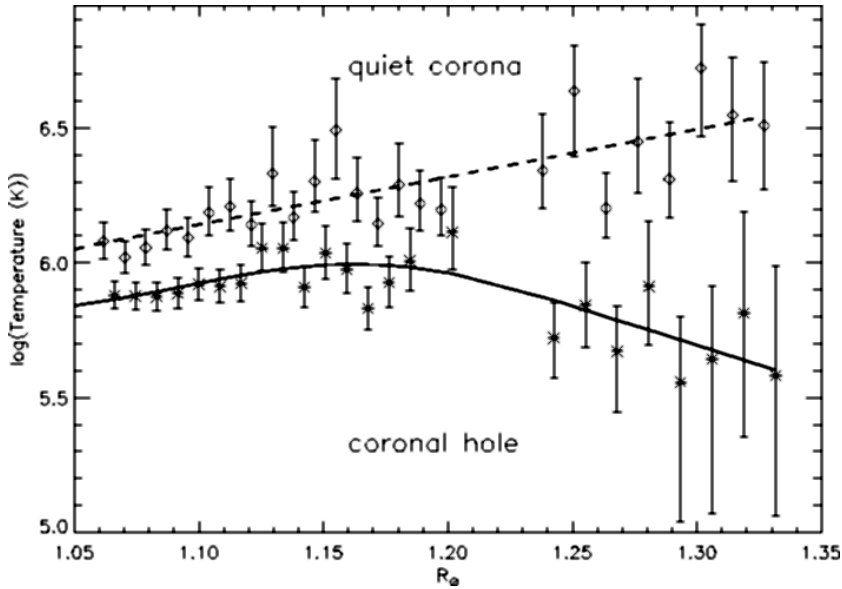


Figure 6(b). Temperature gradient measurement in the quiet corona (equatorial west limb) and the north polar coronal hole (from David *et al.* 1998).

to around 0.4 MK at $1.3 R_{\odot}$. These observations preclude the existence of temperatures over 1 MK at any height near the centre of a coronal hole. Wind acceleration by temperature effects is, therefore, inadequate as an explanation of the high-speed solar wind, and it becomes essential to look for other effects, involving the momentum and the energy of Alfvén waves.

That the solar wind is emanating from coronal holes (open magnetic field regions in the corona) has been widely accepted since the Skylab era. But there was little additional direct observational evidence to support this view. Hassler *et al.* (1999) found the Ne VIII emission blue shifted in the north polar coronal hole along the magnetic network boundary interfaces compared to the average quiet-Sun flow (cf., Fig. 7). These Ne VIII observations reveal the first two-dimensional coronal images showing velocity structure in a coronal hole, and provide strong evidence that coronal holes are indeed the source of the fast solar wind.

Tu *et al.* (2005) have now successfully identified the magnetic structures in the solar corona where the fast solar wind originates. Using images and Doppler maps from the SUMER spectrograph and magnetograms from the MDI instrument on the SOHO

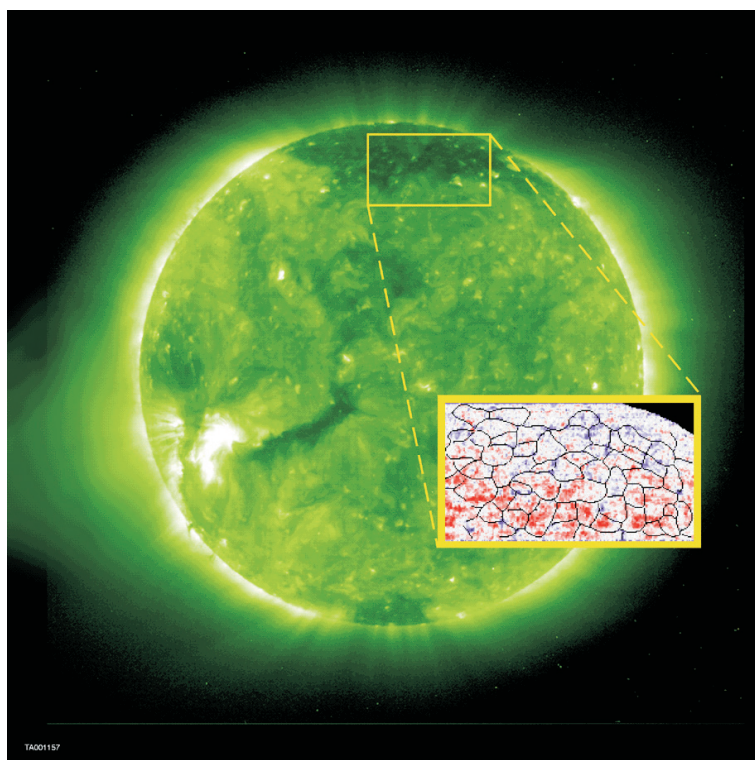


Figure 7. The solar corona and polar coronal holes observed from EIT and SUMER instruments on SOHO. The ‘zoomed-in’ or ‘close-up’ region in the image shows a Doppler velocity map of million degree gas at the base of the corona where the solar wind originates. Blue represents blue shifts or outflows and red represents red shifts or downflows. The blue regions are inside a coronal hole or open magnetic field region, where the high-speed solar wind is accelerated. Superposed are the edges of ‘honeycomb’-shaped patterns of magnetic fields at the surface of the Sun, where the strongest flows (dark blue) occur (Credit: Don Hassler and SUMER-EIT/SOHO).

spacecraft, they have reported the solar wind flowing from funnel-shaped magnetic fields which are anchored in the lanes of the magnetic network near the surface of the Sun. This landmark research leads to a better understanding of the magnetic nature of the solar wind source region. The heavy ions in the coronal source regions emit radiation at certain ultraviolet wavelengths. When they flow towards Earth, the wavelengths of the ultraviolet emission become shorter which can be used to identify the beginning of the solar wind outflow.

Previously it was believed that the fast solar wind originates on any given open field line in the ionization layer of the hydrogen atom slightly above the photosphere. However, the low Doppler shift of an emission line from carbon ions shows that bulk outflow has not yet occurred at a height of 5000 km. The solar wind plasma is now considered to be supplied by plasma stemming from the many small magnetic loops, with only a few thousand kilometers in height, crowding the funnel. Through magnetic reconnection plasma is fed from all the sides to the funnel, where it may be accelerated and finally form the solar wind. The fast solar wind starts to flow out from the top of funnels in coronal holes with a flow speed of about 10 km s^{-1} . This outflow is seen as large patches in Doppler blue shift (hatched areas in the image) of a spectral line emitted by Ne^{+7} ions at a temperature of 600,000 K, which can be used as a good tracer for the hot plasma flow (see Fig. 8). Through a comparison with the magnetic

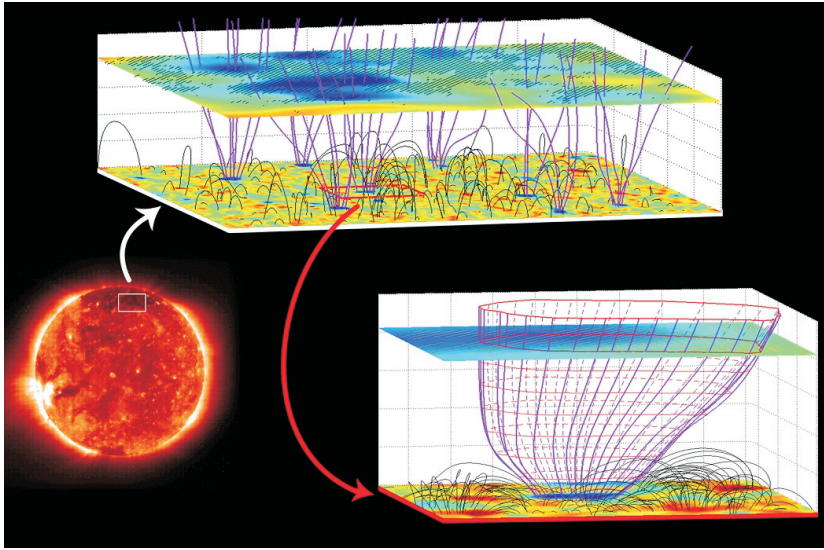


Figure 8. This picture was constructed from measurements which were made on September 21, 1996 on SOHO with the SUMER for Doppler spectroscopy of the coronal plasma, with the MDI for magnetograms of the Sun's surface, and the EIT for the context image of the Sun. The figure illustrates location and geometry of three-dimensional magnetic field structures in the solar atmosphere. The magenta coloured curves illustrate open field lines, and the dark gray solid arches show closed ones. In the lower plane, the magnetic field vertical component obtained at the photosphere by MDI is shown. In the upper plane, inserted at 20,600 km, the Ne VIII Doppler shift is compared with the model field. The shaded area indicates where the outflow speed of highly charged neon ions is larger than 7 km s^{-1} . The scale of the figure is significantly stretched in the vertical direction. The smaller figure in the lower right corner shows a single magnetic funnel, with the same scale in both vertical and horizontal directions (Tu *et al.* 2005) (Credit: SUMER, ESA/NASA).

field, as extrapolated from the photosphere by means of the MDI magnetic data, it has been found that the blue-shift pattern of this line correlates best with the open field structures at 20,000 km (Tu *et al.* 2005).

6. Concluding remarks

A large amount of information concerning the physics of the Sun's hot atmosphere has been made available in recent years from highly successful spacecraft and from ground-based instruments. We have made tremendous progress in pinpointing the processes that maintain the Sun's hot corona and accelerate the solar wind as well as the identification of its source region. And the search goes on.

Acknowledgements

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Magnetic Source Regions of Coronal Mass Ejections

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Abstract. The majority of flare activity arises in active regions which contain sunspots, while Coronal Mass Ejection (CME) activity can also originate from decaying active regions and even so-called quiet solar regions which contain a filament. Two classes of CME, namely flare-related CME events and CMEs associated with filament eruption are well reflected in the evolution of active regions. The presence of significant magnetic stresses in the source region is a necessary condition for CME. In young active regions magnetic stresses are increased mainly by twisted magnetic flux emergence and the resulting magnetic footpoint motions. In old, decayed active regions twist can be redistributed through cancellation events. All the CMEs are, nevertheless, caused by loss of equilibrium of the magnetic structure. With observational examples we show that the association of CME, flare and filament eruption depends on the characteristics of the source regions:

- the strength of the magnetic field, the amount of possible free energy storage,
- the small- and large-scale magnetic topology of the source region as well as its evolution (new flux emergence, photospheric motions, cancelling flux), and
- the mass loading of the configuration (effect of gravity).

These examples are discussed in the framework of theoretical models.

Key words. Coronal mass ejection—flare—filament eruption.

1. Introduction

Coronal mass ejections (CMEs) are large expulsions of mass and magnetic field from the Sun to the interplanetary medium. These mass ejections carry a bulk of material of around 10^{15} g at speeds of 100 to 2700 km s⁻¹ (mean value 487 km s⁻¹ according to Gopalswamy, this issue). In order to forecast CMEs, it is important to understand where they come from on the Sun and how they are related to other flare phenomena (flares, eruptions in particular). With SOHO, it becomes clear that some CMEs are associated with filament eruption occurring in quiet regions of the Sun (decaying active region or polar crown filament). Several statistic analysis have been done to classify CME related to flares, active regions (AR), quiet regions (Saint Cyr & Webb 1991; Webb 1998; Zhou *et al.* 2003). Subramanian & Dere (2001), based on a sample

of 32 CMEs, found that 85% of them are associated with ARs and 15% with so-called quiet regions. Zhou *et al.* (2003) with a sample of 197 front side halo CMES observed by LASCO coronagraphs between 1997 and 2001, find respectively the ratios of 79% and 21% for CMEs associated with AR or not. These two statistics are consistent.

The main characteristics of the source regions have been intensively described in the reviews of Wang (2002) and Schmieder & van Driel-Gesztelyi (2005). CMEs are large scale or global scale activity. An average size of CMEs is as big as five active regions. The CME activity must be related to some other types of large scale magnetic entities. Different types of large scale magnetic structures have been tentatively identified: giant magnetic loops connecting two active regions in opposite hemispheres of the Sun (Delannée and Aulanier 1999) or giant filaments in filament channel (polar crown filament, see Zhou *et al.* 2003) or two rows of opposite polarity field extending to more than 60 heliographic degrees (Wang *et al.* 2002). An example of such a long filament was the transequatorial filament observed before the Bastille Day activity in 2000 (Wang 2002). The destabilization of this huge filament precedes the onset of the flare by many hours. Type II radio storm emissions and radio observations confirm the globalness of the activity. CMEs are also observed after the birth of transient coronal holes (Srivastava *et al.* 2000). These holes of coronal emission are of a different nature than the large dimming regions appearing during the launch of CMEs. Falconer *et al.* (2002) and Deng *et al.* (2001) found correlations of CMEs and global magnetic non-potential configuration. Yan *et al.* (2001) identified the formation of flux ropes in a non-linear force-free-field extrapolation approach.

On the other hand, CMEs have commonly small scale core fields limited to an active region which can be as small as a pore or a new emerging flux (Mandrini *et al.* 2005). Zhang *et al.* (2001) identified magnetic field cancellation associated to CME initiation. In conclusion, the morphology classification of the types of source regions does not bring a good insight of the conditions for getting CMEs. Let us define the necessary conditions for CMEs with and without flares or eruptions.

2. Conditions for CMEs

2.1 General conditions

There is a consensus that the presence of significant magnetic stresses in the source region is a necessary condition for CME.

In the corona the magnetic field can be computed by extrapolation using different approaches: potential, linear force free field or non-linear force-free field. The magnetic helicity quantifies how the magnetic field is sheared or twisted compared to its lowest energy state (potential field). Observations provide plenty of evidence for the existence of such stresses in the solar magnetic field and their association with flares and CME activity.

In young active regions, magnetic stresses are increased by (i) twisted magnetic flux emergence and the resulting magnetic footpoint motions (Wang *et al.* 1994; Leka *et al.* 1996; López-Fuentes *et al.* 2003; Schmieder *et al.* 2002) as well as (ii) torsional Alfvén waves which bring up helicity from the sub-photospheric part of the flux tube and replenish coronal helicity after CME events (Longcope & Welsch 2000; Démoulin *et al.* 2002; Green *et al.* 2002). A possible manifestation of such torsional

Alfvén wave is sunspot rotation, which has indeed been observed (e.g., Brown *et al.* 2003). In old, decayed active regions twist can be redistributed through cancellation events transferring helicity from small towards the large scales, it can be increased by large scale photospheric motions (differential rotation) (DeVore 2000; Démoulin *et al.* 2002b; Berger & Ruzmaikin 2000) or brought up by torsional Alfvén waves (Démoulin *et al.* 2002a; Green *et al.* 2002). Whether the eruption is possible or not, the strength of the magnetic field overlying a sheared arcade or a twisted flux tube may play an important role. Since such an overlying field has a stabilizing effect, a strong field can actually prevent the eruption (Török *et al.* 2004; Roussev *et al.* 2003).

2.2 Current models for the initiation of CMEs

Current models for the initiation of flares, eruptions or CMEs share the assumption that the stored energy lies in a low lying magnetic flux system (twisted or sheared). This system is the so-called core flux or flux rope. The models differ in the way the eruption is initiated and the precise role played by the overlying field which is assumed to be nearly current-free. All these models naturally lead to the configuration of the so-called standard model for the main phase of eruptions (Shibata 1999; Lin & Forbes 2000), in which the long lasting energy release is driven by fast reconnection. The standard model still misses a link between the core flux and the full 3D expansion CMEs. I will not review the CME models but classify the current models of CME's initiation. We can distinguish three types of such models: Forbes & Isenberg (1991) and Isenberg *et al.* (1993) developed 2D and 2.5D models where the eruption is preceded by converging photospheric motions towards the magnetic inversion line of a bipole (Fig. 1). A loss of equilibrium leads to the rise of a flux rope situated along the inversion line. Reconnection occurring in the current sheet formed *under* the flux rope allows the flux rope to escape.

In 3D flux rope models an ideal MHD instability is assumed to destabilize single flux tube anchored in the photosphere. Relevant instabilities are a global expansion of a highly twisted rope (Titov & Démoulin 1999) or the kink instability of a twisted flux tube (Török & Kliem 2003; Török *et al.* 2004). The emerging twisted flux rope model is called tether cutting model in the classification of storage and release energy models of Klimchuk (2001) (Fig. 2). It implies magnetic reconnection *within* the core flux. Finally the third model is the breakout model (Antiochos 1998; Antiochos *et al.* 1999; Aulanier *et al.* 2000; Wang 2002) where magnetic reconnection is assumed to be *above* the core flux and breaks its balance with the overlying flux, thus initiating an eruption. A quadrupolar magnetic configuration is required in this model.

I shall review a few examples and point out the main important factors which could be responsible for the CME and what type of models would be most relevant to explain the CME.

- High field strength and magnetic stress
- Evolution of active regions
- Amount of twist: large sigmoids
- Mass loading in prominences

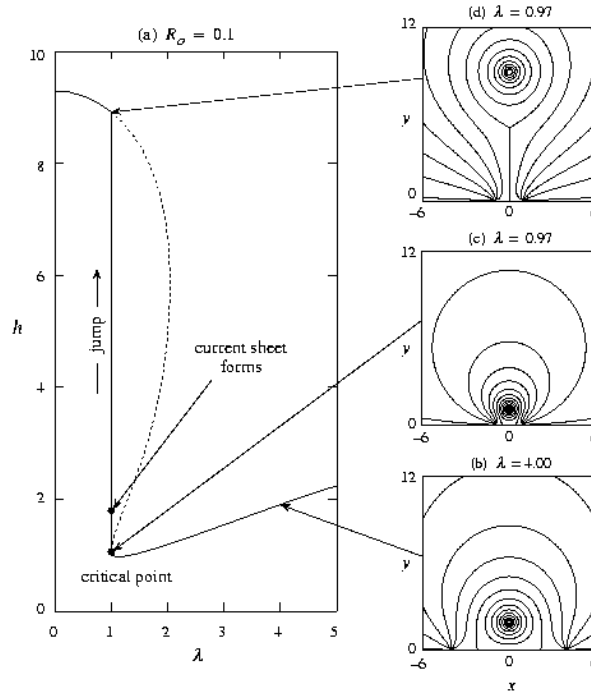


Figure 1. Loss of equilibrium model: sketches of coronal field lines showing converging motions of the footpoints of overlying arcades leading to loss of equilibrium of the structure (Isenberg *et al.* 1993). Reconnection occurs below the magnetic flux rope.

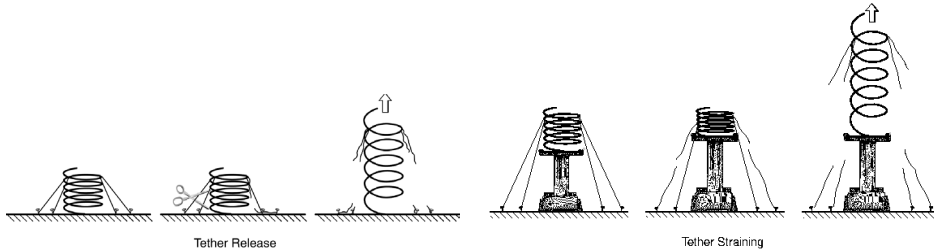


Figure 2. Sketches of models of energy release, left: tether cutting, right: tether straining (Klimchuk 2001).

3. Review of examples

3.1 Large strength of magnetic field and high stress

Commonly, active regions with strong magnetic field and high stresses are sites of large energetic flares producing fast CMEs (e.g., the Bastille day flare, the Halloween flares during the period of October–November 2003). The twelve X-ray class Halloween flares come from three complex active regions NOAA 10484, 10486 and 10488. These three source regions are delta-spot regions formed through numerous episodes of flux emergence. As an example, the active region NOAA 10486 has an overall quadrupolar

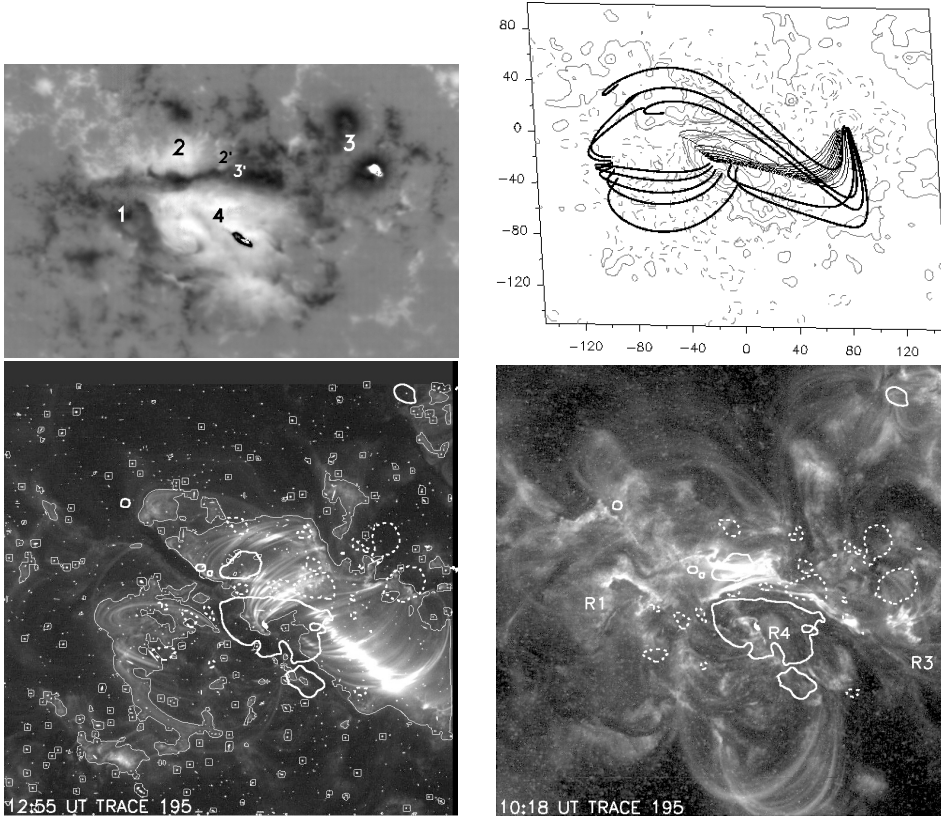


Figure 3. X-ray 17 Flare of October 28, 2003, (**top left**) magnetogram of the active region 10486 observed with THEMIS showing the four main polarities, (**top right**): magnetic field lines of the quadrupolar reconnection before the flare, (**bottom left**): TRACE observations of the loops after the X17 flare, (**bottom right**) TRACE observations of the four ribbons signature of the quadrupolar reconnection before the flare (Schmieder *et al.* 2006).

magnetic configuration on October 28, 2003 (Schmieder *et al.* 2006; Mandrini *et al.* 2006). The magnetic field strength was still high although the active region had entered into its decaying phase, when an X17 GOES class flare and associated CME occurred (Fig. 3). The magnetograms of Huairou indicate continuous emergence of new flux and the build-up of a strong shear along the two parallel inversion lines which squeezed a bridge of negative polarities between two positive polarities (Zhang *et al.* 2003).

3.2 Evolution of active regions

Comprehensive analyses of the long term evolution of two active regions confirm that evolving magnetic flux density plays an important role in CME activity and its evolution must influence all the activity signatures (van Driel-Gesztelyi *et al.* 2003; Green *et al.* 2002). The first analysis concerns an isolated active region (NOAA 7978 July–December 1996). The magnetic field of the AR was clearly distinguishable for at least seven solar rotations. It was found that flares mainly occur when the magnetic field of the AR has the highest complexity and magnetic flux density during the two

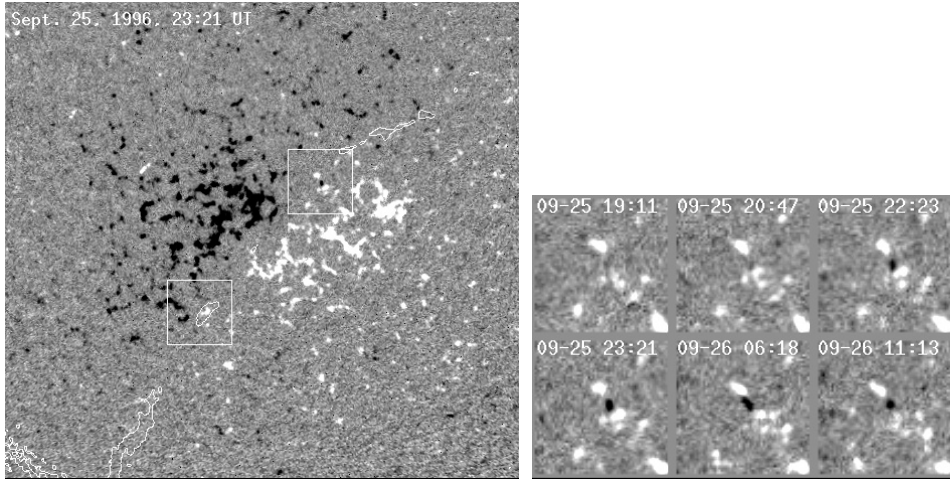


Figure 4. (Left panel) SOHO/MDI magnetogram of September 25, 1996 at 23:21 UT with the 2 boxes where we follow the evolution of the small polarities. The contours indicate the position of the H α filament. (Right panel) New emergence of magnetic field and cancellation observed versus time in the upper box (Schmieder *et al.* 2002).

main flux emergence phases (1st two rotations in July and August 1996) and while the number of CMEs is sustained and may be more closely related to the magnetic stresses in the region, since the value of the linear force-free α parameter was found to be roughly proportional with the CME rate. During the first observed flux emergence phase 10^{22} Mx flux surfaced. During the decay phase magnetic flux gradually spreads over an ever increasing area, flare activity shows a sharp decrease and practically ceases with the disappearance of sunspots, while CME activity remains on a relatively high level (3–4 CMEs per rotation). During the decaying phase of the active region **cancelling flux** along the major inversion line was identified and could participate to the redistribution of the twist (Fig. 4). The filament which was lying along the inversion line erupted after a sustained period of cancellations (Schmieder *et al.* 2002). The position and the time of this flux cancellation lead us to propose that it could be responsible for the destabilization of the filament, which was probably close to loss equilibrium already. A similar evolutionary pattern was found in NOAA AR 8100 by Green *et al.* (2002).

3.3 Sigmoids of small and large scale

Sigmoids, in many cases, indicate the presence of high magnetic stresses and have been linked to CMEs (Gibson *et al.* 2002; Manoharan *et al.* 1996). Though the sigmoid–CME connection is still statistically ambiguous (Canfield *et al.* 1999; Aulanier *et al.* 2005) it has been suggested that the magnetic helicity content in S-shaped magnetic configuration may reach a threshold leading to instability and eruption (van Driel-Gesztelyi *et al.* 2000; Török & Kliem 2003). The sigmoid observed on 25 Oct. 1994 by Yokoh is a good example to understand the process leading to CME (Manoharan *et al.* 1996). In soft X-rays the region had a sigmoidal shape (Fig. 5). Van Driel-Gesztelyi *et al.* (2000) proposed a model involving reconnection in a sheared arcade, which

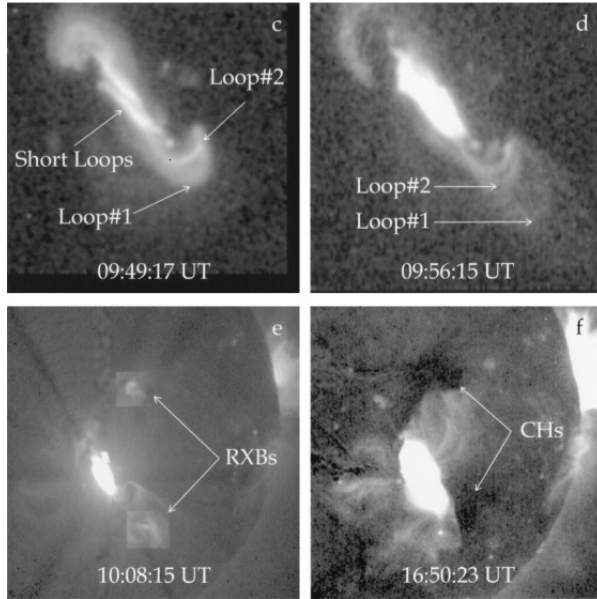


Figure 5. Sigmoid in X-ray before the flare and the reconnected loops after the flare observed with Yohkoh (Manoharan *et al.* 1996).

leads to the formation of short, highly sheared loops in the centre and long sigmoidal loops connecting the outer edges of the bipole. The long sigmoidal loops may become unstable leading to a CME. As the sigmoid expands, a current sheet is formed under it and a cusp structure appears.

A similar event was observed in the center of the disk but related to a very small bipole, which had a magnetic flux which was two orders of magnitude less than a usual active region. Mandrini *et al.* (2005) provided several independent supporting arguments for the eruptive nature of this tiny active region, linking the sigmoid eruption to a small interplanetary magnetic cloud.

4. Role of magnetic topology

4.1 Bipolar or quadrupolar magnetic configuration

The coronal environment of the ejection of coronal mass is certainly important to produce fast and large heliographic extended CMEs (Low & Zhang 2002). In this review I shall not consider the global structure of the CME source environment. The environment defines mainly the morphology of the CME but not the reason of the launch of a CME. I shall concentrate on the photospheric magnetic configuration of sources. Do CME source regions have bipolar or quadrupolar magnetic structure? That is not clear.

In bipolar region, converging motion of opposite polarities and consequent cancellation along the magnetic inversion line before CMEs have indeed been observed (van Driel *et al.* 2002). It was the case for the CME of September 1996 before the eruption of the filament as we have already shown (Fig. 4). This latter example is relevant to

the loss equilibrium proposed by Forbes & Isenberg (1991) in a bipolar region. Such an eruption could also be relevant to the tether cutting model. The filament starts to rise slowly, as more and more of the tethers have been cut. During that time the twist in the flux tube increases.

However, it is still a question of whether or not topological complexity is indispensable for CMEs. Does the loss of equilibrium of a flux rope, in which twist exceeded threshold level (see Török *et al.* 2004 and references within), provide sufficient condition for CME? The latter seems to work even in a simple bipolar magnetic configuration.

However, the highly CME-active source regions are magnetically complex! What is the main role of a multipolar magnetic structure in the CME process?

From the analysis of the magnetic field evolution, topology and multi-wavelength data before and during the X 17 flare on October 28, 2003, we conclude that a large twisted flux tube supporting the long filament lying along the main inversion line must have been formed and erupted. The slow build-up of magnetic stress through converging motions along the magnetic inversion line begins well before the eruption. If we try to classify the X 17 flare within the scenario of ‘storage and release’ models (Klimchuk 2001) i.e., tether cutting or tether straining (Fig. 2), it appears that two mechanisms could be present: firstly, the twist in the flux rope built up as small-scale cancellation events (reconnection) occurs in a sheared arcade aligned along the main neutral line. The flux tube starts to rise slowly, as more and more tethers are being cut. During that time the twist in the tube increases. Before the X 17 flare, two episodes of large-scale quadrupolar reconnections occur in the active region (Schmieder *et al.* 2006). The second episode, being more intense, implies more important field-rearrangements. These quadrupolar reconnection events remove stabilising field lines from above the flux rope (filament) which succeeds to break out. Reconnection under the erupting flux rope starts after the lift-off, forming a post-flare loop arcade. For this flare it appears that first there is reconnection in the low atmosphere. The stored energy could only escape when the overlying field lines reconnected in the higher corona. This evolution is similar to the breakout model. Similar scenarios were proposed for the 14 July 1998 CME by Aulanier *et al.* (2000) and for the 15 July 2002 CME by Gary & Moore (2004). In the breakout models the quadrupolar magnetic configuration is a necessary condition.

4.2 Magnetic field strength

It has been a long-standing question regarding which comes first: flare or CME? Both observations and models suggest that an inflation of the magnetic structure due to increasing magnetic stresses is the first step – this makes observers say that CME starts before flare. The main flare energy release occurs during the reconnection of field lines extended by the eruption, which, again, points towards that CME should come first. However, both in the tether cutting (flux rope) and the tether straining (breakout) models pre-eruption reconnection is required under or above, respectively, of the erupting twisted/sheared magnetic structure (filament). In the breakout model this pre-eruption reconnection is expected to occur high in the corona in a region of weak field, releasing too little energy to be observed. However, if the reconnection occurs in a strong field region, like in a complex active region, where tethers are not external weak fields overlying the sheared core field but are part of the active region having a quadrupolar configuration, the released magnetic energy can be high enough

to qualify as flare, therefore we may find that an impulsive, quadrupolar flare precedes the CME, while the latter includes a filament eruption and a related two-ribbon flare representing the post-eruption arcade.

In any case, in the models most relevant to observations (storage and release models) the build-up of magnetic stresses, i.e., strong shear and twist are necessary conditions for CME to occur. In observations twisted CME structures are frequently seen and there is a mounting evidence that considerable amount of twist is being carried away from the Sun by CMEs. These emphasize the importance of magnetic helicity in the CME process.

5. Mass loading in prominences

Finally, another important parameter for getting CMEs is the mass loading. The pressure increases over the stressed magnetic field the kinetic energy is increasing during the disruption of the field. This can be the case for filament eruption in or out of active regions. The free energy stored in a stressed magnetic structure prior to the eruption depends on the strength of the background field. The stronger the background field is, the more free energy can be stored, and thus the more energetic the eruptive process will be (Lin 2004). This eruptive process refers to any disruption of the coronal magnetic field that causes either a flare, or a filament eruption or CME or all of them. In the case of CMEs related to quiet sun region, there is a critical strength of the background magnetic field (< 27 G) where the effect of gravity becomes significant enough to prevent the CME from progressing (Isenberg *et al.* 1993). If the field is larger than this value, the system evolves smoothly in response to the slow change in the boundary conditions and can end up with a slow CME (Schmieder *et al.* 2000). The gravity of the filament or prominence may play an important role in the process of slow CMEs by prohibiting the catastrophe to occur at the very beginning. A prominence rises slowly before erupting and can reach an altitude as high as 200 mm. The mass of CMEs inferred from SOLWIND observations from 1979 to 1981 (Howard *et al.* 1985) ranged from 2×10^{14} to 4×10^{16} g. The amount of mass and the mass concentration should be important for the CME-eruptive prominence association. Recently it has been found that solar filaments observed in EUV lines are much more extended than their $H\alpha$ counterparts (Heinzel *et al.* 2001; Schmieder *et al.* 2003). The total filament mass could be larger than that derived for the $H\alpha$ filament itself by a factor 1.5 or 2 and this may have consequences for the structure and the mass loading of CMEs (Heinzel *et al.* 2003).

6. Conclusion

The driver behind solar flares, prominence eruptions and CMEs is the instability of stressed coronal magnetic fields which have been driven towards a highly stressed state by photospheric motions and twisted flux emergence. From a theoretical point of view (Lin 2004) and observations of CME source regions, it appears that some parameters are important to determine which type of CME the region is able to expel. The strength of the magnetic field and the complexity of the magnetic configuration prior to the eruption determines the correlation between flares and CMEs, and the impact of the gravity on the above correlation can be important (mass and concentration of mass

in prominence) if the background magnetic field is weak. The size of the region is not an important parameter, even very small magnetic regions are able to produce CME. An increase of stress in the source region is a crucial factor, and the strength of the field determines how much free energy can be stored in it. However, the strength of the overlying field, which has an important stabilizing effect, appears to play an important role in whether or not eruption occurs. The correlation between CME and eruptive prominence depends on the amount and the concentration of the plasma mass in the related magnetic configuration. The CME may commence with an apparent prominence eruption if the mass inside the structure is larger than 4.5×10^{15} g.

Acknowledgements

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North–South Distribution of Solar Flares during Cycle 23

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Abstract. In this paper, we investigate the spatial distribution of solar flares in the northern and southern hemispheres of the Sun that occurred during the period 1996 to 2003. This period of investigation includes the ascending phase, the maximum and part of the descending phase of solar cycle 23. It is revealed that the flare activity during this cycle is low compared to the previous solar cycle, indicating the violation of Gnevyshev–Ohl rule. The distribution of flares with respect to heliographic latitudes shows a significant asymmetry between northern and southern hemisphere which is maximum during the minimum phase of the solar cycle. The present study indicates that the activity dominates the northern hemisphere in general during the rising phase of the cycle (1997–2000). The dominance of northern hemisphere shifted towards the southern hemisphere after the solar maximum in 2000 and remained there in the successive years. Although the annual variations in the asymmetry time series during cycle 23 are quite different from cycle 22, they are comparable to cycle 21.

Key words. Sun: activity—flares—north–south asymmetry.

1. Introduction

The distribution of various solar activity features with respect to heliographic latitudes as a function of time has been investigated in several studies. These activity features include flares, filaments, magnetic flux, sunspot numbers, sunspot area, etc. These studies indicate that a solar cycle is not symmetric considering the distribution of solar activity separately in northern and southern hemisphere. This intrinsic feature (N–S asymmetry) poses a challenge for dynamo model calculations.

Howard (1974) examined the N–S distribution of solar magnetic flux for the period 1967–1973 and concluded that about 95% of the total magnetic flux of the Sun is confined to latitudes below 40° in both the hemispheres. It was also found that total magnetic flux in the north exceeded to that in the south by 7%. Roy (1977) studied the N–S distribution in the data of major flares, sunspot area and their magnetic configuration during 1955–1974 and found the dominance of northern hemisphere over the southern one in all these categories. Garcia (1990) investigated the N–S distribution of soft X-ray flares (class $\geq$ M1) during solar cycle 20 and 21. It was concluded that the spatial distribution of flares varies within a solar cycle such that the preponderance of flares occurs in the north during the early part of the cycle and then moves south

as the cycle progresses. Knořka (1985) investigated the N–S asymmetry of H α flare index, introduced by Kleczek (1953), for cycles 17–20, which was later extended till cycle 22 by Ata & Özgü (2001). Their results show a long-term periodic behaviour in asymmetry time series. Verma (1993) examined the N–S asymmetry of various solar active phenomena and reported cyclic behaviour of asymmetry. Joshi (1995) and Li *et al.* (1998) studied the N–S asymmetry of H α and soft X-ray flares respectively during solar cycle 22. In both the above investigations, a southern dominance was prevalent. Joshi & Pant (2005) analysed the data of solar H α flares during solar cycle 23 to investigate their spatial distribution. Recently Knaack *et al.* (2004, 2005) studied the temporal and spatial variations in the photospheric magnetic flux between the northern and southern hemisphere of the Sun from 1975 to 2003 and reported significant periodic variations of magnetic activity between the two hemispheres. They also studied N–S asymmetry using the monthly averaged sunspot areas obtained for the period of 1874–2003.

In the present study, we investigate the latitudinal distribution of H α flares during the present solar cycle 23. We have studied the yearly variations in N–S asymmetry and discussed the significance of observed asymmetry by using the binomial probability test. These results for solar cycle 23 have been compared with the behaviour of previous cycles 22 and 21 as reported by Temmer *et al.* (2001).

2. H α flare data

The data used in the present study have been collected from H α flare lists published in the SGD (Solar Geophysical Data) during the time span of 01 May 1996 to 31 December 2003, covering almost 8 years of solar cycle 23. During this period, the occurrence of 20235 H α flares is reported. In H α , flares are classified according to their importance and brightness classes. The important class (S = subflare, or 1, 2, 3, or 4 for successively large flares) denotes the size of the flare and the brightness class (f = faint, n = normal, b = bright) corresponds to a subjective estimate of the intensity of the emission. In the list of H α flares, classifications are not given for some events. After excluding such events, we get a total of 20223 H α flares, which provides an extensive database for the present study. Figure 1 shows the plot of monthly flare counts as well as monthly mean sunspot numbers during the period of our investigation.

3. Latitudinal distribution and N–S asymmetry

To study the spatial distribution of flares with respect to heliographic latitudes, we have calculated the number of flares in the interval of 10° latitude for northern and southern hemispheres (see Table 1). In this table those events have been excluded which occurred at 0° latitude. Since the number of flares above 50° latitude is very small in both the hemispheres, the number of flares occurring above 50° latitude is merged in one group. Column 8 of Table 1 gives the total number of flares in the northern and southern hemispheres. It is evident from Table 1 that for all these years one hemisphere produces more flares than the other.

It is customary to describe N–S asymmetry by an asymmetry index

$$A = \frac{N - S}{N + S}, \quad (1)$$

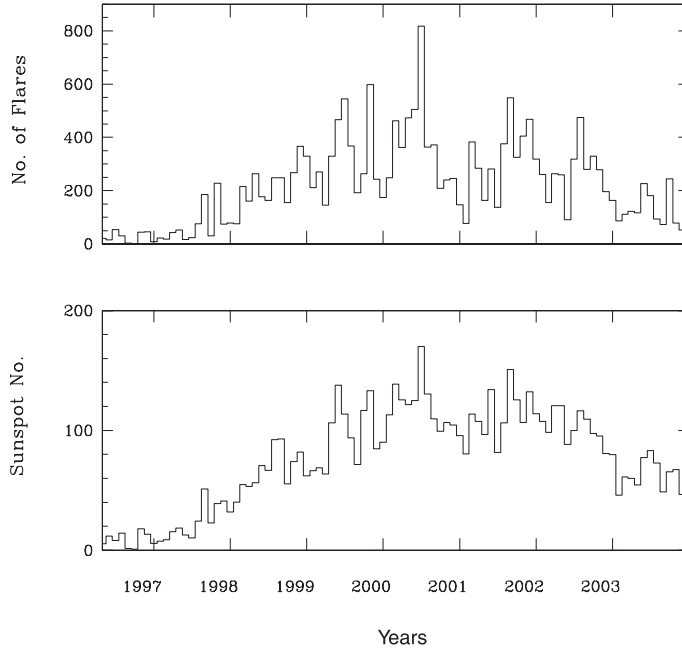


Figure 1. Monthly numbers of H α flares and monthly mean sunspot numbers from 1996–2003.

where N and S are the yearly number of flares in the northern and southern hemisphere of the Sun respectively. The asymmetry indices, based on annual flare counts from 1996 to 2003, have been plotted in Fig. 2. The statistical significance of the flare dominance in northern and southern hemispheres has been assessed by using the binomial probability distribution. Let us consider a distribution of n objects in 2 classes. The binomial formula gives us the probability $P(k)$ of getting k objects in class 1 and $(n - k)$ objects in class 2, such that

$$P(k) = \frac{n!}{k!(n - k)!} p^k (1 - p)^{n-k} \quad (2)$$

and the probability to get more than d objects in class 1 is given by

$$P(\geq d) = \sum_{k=d}^n P(k). \quad (3)$$

In general, when $P(\geq d) > 10\%$, implies a statistically insignificant result (flare activity should be regarded as being equivalent for the two hemispheres), when $5\% < P(\geq d) < 10\%$ it is marginally significant, and when $P(\geq d) < 5\%$ we have a statistically significant result (flare occurrence is not due to random fluctuations). The calculated values of probability are given in Table 1 and based on that the highly significant, marginally significant and insignificant values of asymmetry indices are marked with different symbols in Fig. 2. Figure 2 also shows the yearly N–S asymmetry (given by equation 1) of H α flare index (dashed line) from 1996 to 2003.

Table 1. Number of H α flares at different latitude bands in the northern (N) and southern (S) hemispheres are tabulated for each year. The binomial probability (Prob.) and the dominant hemisphere (DH) is given for all the years as well as for all the latitudinal bands. Dash (–) represents that the probability is not significant. Flares occurred exactly at the equator have been excluded.

Years	Number of flares						Total	Prob.	DH
	0–10°	10–20°	20–30°	30–40°	40–50°	> 50°			
1996 N	21	2	0	1	0	0	24	1.107×10^{-33}	S
S	112	53	21	2	0	0	188		
1997 N	39	160	230	20	0	0	449	5.641×10^{-6}	N
S	5	103	199	18	1	1	327		
1998 N	13	635	503	54	1	1	1207	0.428	–
S	3	366	807	31	9	0	1216		
1999 N	169	1240	811	105	11	0	2336	7.017×10^{-30}	N
S	71	874	644	38	0	0	1627		
2000 N	463	1327	633	55	1	2	2481	8.712×10^{-14}	N
S	248	1288	373	79	1	0	1989		
2001 N	503	978	262	5	1	0	1749	0.058	S
S	449	1062	282	42	8	0	1843		
2002 N	250	746	255	6	0	0	1257	1.329×10^{-35}	S
S	675	1006	272	4	0	0	1957		
2003 N	407	315	14	3	0	0	739	0.069	S
S	222	475	83	17	0	0	797		
Total N	1865	5403	2708	249	14	3	10242	0.018	N
S	1785	5227	2681	231	19	1	9944		
Prob. DH	0.093 N	0.956 –	0.356 –	0.205 –	0.189 –	0.125 –	0.018 N		

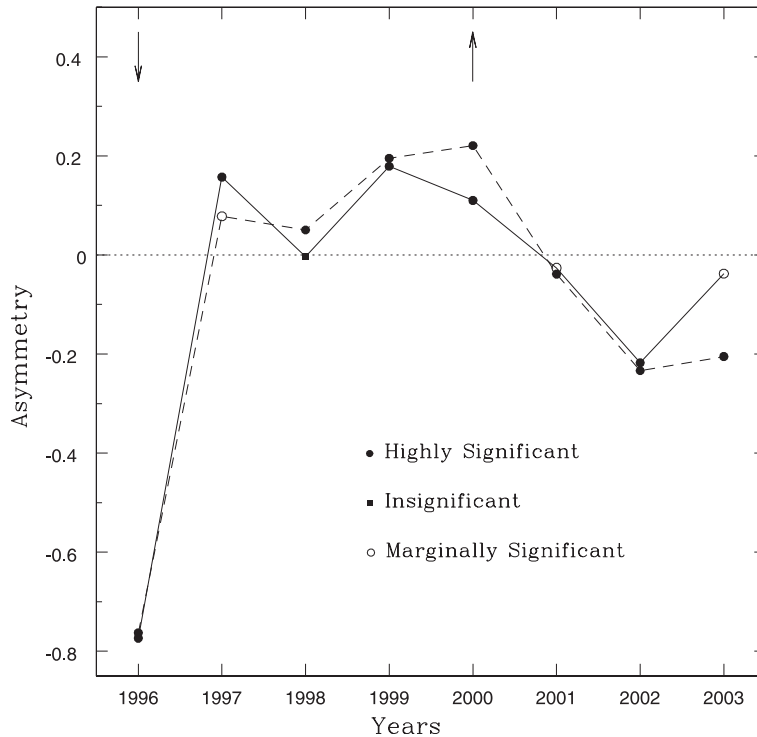


Figure 2. Plot of annual N–S asymmetry index of H α flare counts (solid line) and H α flare index (dashed line) from 1996 to 2003. Arrows indicate the maximum and minimum phases, respectively, of solar cycle 23.

4. Discussions and conclusions

Temmer *et al.* (2001) and Temmer (2004) made an extensive statistical analysis of H α flare data of cycles 21 and 22 to study their temporal behaviour and spatial distribution over the solar cycle. The comparison of monthly flare counts during cycle 23 (Fig. 1) with those reported by Temmer *et al.* (2001) for cycles 22 and 21 shows that the activity level during cycle 23 is significantly lower than two previous cycles. This clearly indicates the violation of the Gnevyshev–Ohl (G–O) rule in terms of the level of flare activity, for the pair of solar cycles 22–23. This empirical rule states that the sum of sunspot numbers over an odd cycle exceeds that of the preceding even cycle (Gnevyshev & Ohl 1948). The sunspot numbers, starting from cycle 0 (i.e., from the year 1750), show that G–O breakdown had also occurred for the Hale cycles consisting of the 11-year pair of cycles 4–5 and 8–9. Komitov & Bonev (2001) examined the conditions for the violation of the G–O rule. They analysed a long data set of 152 solar activity cycles obtained from direct and indirect records and predicted a high probability for violation of the G–O rule for the pair of the cycles 22–23.

Table 1 shows several interesting aspects of flare distribution with the evolution of solar cycle. In the beginning of the cycle 0–10° latitudinal belt in southern hemisphere produced maximum number of flares, which could be a remnant of the preceding cycle. In the year 1997, just after solar minimum, most of the flares were produced in

the 20–30° latitudinal belt and with the progress of solar cycle, the flare occurrence increased in lower latitudes also. Table also shows that 10–20° latitudinal belt was the highest flare producing region.

Figure 2 shows the annual variations in N–S asymmetry. There is a strong southern dominance during the solar cycle minima in 1996. This behaviour of asymmetry that it peaks at or around the minimum phase of solar activity has been reported in several studies with different manifestations of solar activity (Swinson *et al.* 1986; Vizoso & Ballester 1990; Joshi & Joshi 2004). In 1997, 1999 and 2000, when the cycle was in the ascending phase, northern hemisphere dominated. The preference for northern hemisphere during the rising and maximum phase of cycle 23 is reported by Ataç & Özgüç (2001) and Joshi & Joshi (2004) in yearly values of $H\alpha$ and soft X-ray flare index respectively. Gopalswamy *et al.* (2003) compared the latitudinal distribution of prominence eruptions (PEs) and coronal mass ejections (CMEs) as a function of time from 1996 to 2002. Their study revealed that there is a shift in the dominance of PEs and CMEs activity from northern to southern hemisphere after solar maximum in 2000. Similar trend has been found in the present investigation with solar flare count data. The yearly variations in N–S asymmetry for flare counts and flare index (Fig. 2) show a similar trend. In the year 2000 flare index asymmetry is stronger in the northern hemisphere compared to flare count asymmetry, which indicates that most of the major flares occurred in the northern part of the solar disk during solar cycle maxima. Similarly during October and November months of 2003, most of the big flares occurred in the southern hemisphere of the Sun and consequently the asymmetry in flare index values became stronger in the southern part of solar disk. Comparing our results with Temmer *et al.* (2001) we find that the variations in N–S asymmetry index during cycle 23 differ from cycle 22 but are similar to cycle 21. There was mostly a southern dominance during cycle 22 while cycle 21 showed northern dominance during the early phases and southern dominance during the later phases. The present investigation is consistent with the N–S asymmetry analysis performed with soft X-ray flare index data during cycles 21–23 (Joshi & Joshi 2004).

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Initiation of CMEs by Magnetic Flux Emergence

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Abstract. The initiation of solar Coronal Mass Ejections (CMEs) is studied in the framework of numerical magnetohydrodynamics (MHD). The initial CME model includes a magnetic flux rope in spherical, axisymmetric geometry. The initial configuration consists of a magnetic flux rope embedded in a gravitationally stratified solar atmosphere with a background dipole magnetic field. The flux rope is in equilibrium due to an image current below the photosphere. An emerging flux triggering mechanism is used to make this equilibrium system unstable. When the magnetic flux emerges within the filament below the flux rope, this results in a catastrophic behavior similar to previous models. As a result, the flux rope rises and a current sheet forms below it. It is shown that the magnetic reconnection in the current sheet below the flux rope in combination with the outward curvature forces results in a fast ejection of the flux rope as observed for solar CMEs. We have done a parametric study of the emerging flux rate.

Key words. Sun: corona, coronal mass ejections—magnetic fields—magnetohydrodynamics—flux emergence.

1. Introduction

Coronal Mass Ejections (CMEs) play a *crucial role* in space weather and a careful study of the origin, structure, and propagation characteristics of this violent phenomenon is essential for a deeper insight into space weather physics. The fast CMEs are important because of the geomagnetic storms created by the impact of the CMEs on the earth's magnetosphere and the solar energetic particles (SEPs) accelerated by the CME leading shock front. Clearly, the theoretical modelling of the evolution of CMEs can be divided into sub-problems. An important unresolved sub-problem is the initiation of CMEs, i.e., the question why CMEs occur at all and how are they triggered?

Klimchuk (2001) reviewed the theoretical models for CME initiation. Based on basic physical properties such as energetics, structure and dynamics, this author distinguished two types of models, viz., *directly driven* models and *storage and release* models. However, *all these models have difficulty explaining one or more aspects of observations*. Hence, a lot of work remains to be done to improve the present models: 3D extensions need to be created including fine structure, real (i.s.o. numerical) dissipation, a realistic shearing of foot points, etc. CME shock evolution studies very often

apply *simplified* generation models for CMEs. Groth *et al.* (2000) used a ‘density-driven’ model which simply means that a high-density (and high-pressure) plasma blob is superposed on the solar wind. Other simplified models generate a pressure pulse with or without an additional velocity change (i.e., a kind of ‘nozzle’ boundary condition), see e.g., Wang *et al.* (1995) and Odstrcil & Pizzo (1999). Keppens & Goedbloed (1999), on the other hand, impose an extra mass flow to generate CMEs.

More realistic CME evolution simulations make use of theoretical analytic CME initiation models such as the models of the self-similar Gibson & Low (1998) family (Gombosi *et al.* (2000)) or the Titov & Démoulin model (Roussev *et al.* (2003)). Odstrcil & Pizzo (2002) and Odstrcil *et al.* (2003) use yet another analytical flux rope model. What one should really do, however, is to simulate the evolution of the reconstructed coronal structures (e.g., Aulanier *et al.* (2000)) driven unstable by foot point shearing and/or flux emergence or cancellation. Several groups are working on such simulations and results will appear in the near future.

The CME initiation model presented in this paper extends earlier results obtained by Lin *et al.* (1998) and Chen & Shibata (2000) by including geometry effects (curvature) and gravitational stratification, and performing a parametric study.

2. Model and numerical method

2.1 Initial model

Our pre-CME model is motivated by the work of Chen & Shibata (2000) and Lin *et al.* (1998). Chen & Shibata used three separate line current elements to get the detached magnetic flux rope in a 2D cartesian plane. We have also considered three separate current elements but in a 2.5D spherical plane: a ring current centered at distance h above the photosphere on the equator with a finite radius r_0 , its image current is located below the photosphere, and a background dipole field, similar to that in Forbes (1990). In the present work, we do not consider the formation of flux ropes but in previous studies, several mechanism have been suggested for this.

We use spherical coordinates (r, θ, φ) and express the magnetic field $\mathbf{B}$ in terms of a vector potential A_φ , i.e., $\mathbf{B} = \nabla \times (A_\varphi \mathbf{e}_\varphi) + B_\varphi \mathbf{e}_\varphi$. The initial magnetic configuration is given by $A_\varphi = A_{\varphi r} + A_{\varphi i} + A_{\varphi b}$, where the magnetic vector potential of the ring current ($A_{\varphi r}$), its image current ($A_{\varphi i}$) and the background dipole field ($A_{\varphi b}$) have the following forms:

$$A_{\varphi b}(r, \theta) = m \frac{\sin \theta}{r^2}, \quad (1)$$

$$A_{\varphi r}(r, \theta) = \begin{cases} \frac{4Ih}{\sqrt{h^2 + r^2 + 2hr \sin \theta}} \frac{[(2 - k_1^2)K(k_1) - 2E(k_1)]}{k_1^2} & r_1 > r_0 \\ B_0 \frac{j_0(\lambda r_1)}{\lambda r \sin \theta} & r_1 \leq r_0, \end{cases} \quad (2)$$

$$A_{\varphi i}(r, \theta) = -\frac{4IhR_0}{\sqrt{r^2 h^2 + R_0^4 + 2hrR_0^2 \sin \theta}} \frac{[(2 - k_2^2)K(k_2) - 2E(k_2)]}{k_2^2}, \quad (3)$$

where m is the dipole strength in the formula of background field, j_0 is a Bessel function of zeroth order, $r_1 = \sqrt{r^2 + h^2 - 2rh \sin \theta}$ is the radial distance from the

flux rope center, K and E are the complete elliptic integrals of the first and second kind, respectively, and

$$k_1^2 = \frac{4hr \sin \theta}{h^2 + r^2 + 2hr \sin \theta}, \quad k_2^2 = \frac{4hr R_0^2 \sin \theta}{h^2 r^2 + R_0^4 + 2hr R_0^2 \sin \theta}.$$

The constant λ is chosen such that $j_0(\lambda r_0 = 0)$. Moreover, we consider a flux rope with radius $r_0 = 0.1 R_\odot$ at a distance $h = 1.3 R_\odot$ from the solar center (i.e., $0.3 R_\odot$ above the photosphere). The strength of the magnetic field in the center of the flux rope is chosen as $B_0 = 5.6$ G.

We assume axial symmetry in our model and that there are no magnetic field sources at infinity. The current within the flux rope, that is located at the equatorial plane, is the only current source within the region $r > R_\odot$ prior to the eruption.

We have taken into account the effect of gravity in our model. As a result, the ambient density is stratified, i.e., $\rho_g = \exp((v_{\text{esc}}^2/r) - 1)/2)$, where v_{esc} is the escape velocity of the Sun. The flux rope is initialized with a high density that is distributed as:

$$\frac{\rho}{\rho_0} = \begin{cases} \rho_g + \frac{2 \cos \pi r_1^2}{\beta (2r_0)^2} & r_1 \leq r_0 \\ \rho_g & r_1 > r_0. \end{cases} \quad (4)$$

An initial uniform temperature is assumed $T = T_0$ and the initial velocity $\mathbf{v}$ is set to zero. The initial distribution of the electric current density and magnetic field inside the flux rope is arbitrary to some extent. The initial state is iteratively relaxed to a numerically satisfied equilibrium state before the flux triggering mechanism is applied.

2.2 Numerical methods

The 2.5D (axially symmetric) time dependent ideal MHD equations are applied. This set of PDEs is solved numerically by means of the VAC Code (Versatile Advection Code, Toth (1996)) by a two-step Runge-Kutta scheme in time and for the spatial discretization we use a second order finite volume scheme, viz., the TVD Lax-Friedrichs method. The solenoidal condition, $\nabla \cdot \mathbf{B} = 0$, is maintained to machine precision using Balsara's (2003) approach.

We have taken the following characteristic values for normalizing the density and temperature: $\rho_0 = 1.6726 \times 10^{-16}$ g/cm³, and $T_0 = 1.5 \times 10^6$ K, respectively. The solar radius $R_\odot = 6.9626 \times 10^{10}$ cm is used to normalize the length scales. The numerical units for the other quantities follow from these, viz., velocity $v_0 = T_0 2k_b/m_p = 15.74 \times 10^6$ cm/s, time $t_0 = R_\odot/v_0 = 4420$ s, $B_0 = v_0 \sqrt{4\pi\rho_0} = 0.7216$ Gauss, and magnetic flux $\psi_0 = B_0 R_\odot^2 = 34.98 \times 10^{20}$ Mx. We have considered ideal MHD but reconnections of magnetic field lines can occur due to the numerical diffusion. In the near future, we will take into account finite magnetic resistivity in our model.

The computational domain is part of the (r, θ) -plane limited by $1 \leq r \leq 30$ and $0 \leq \theta \leq \pi$. This domain is discretized by 600×201 grid points, which are non-uniformly distributed in both the r - and θ -directions. The numerical grid is accumulation near $r = 1$ and $\theta = \pi/2$ with a stretch factor of 50 and 10 in the r - and θ -directions, respectively. The bottom boundary of the simulation area is a line tying boundary,

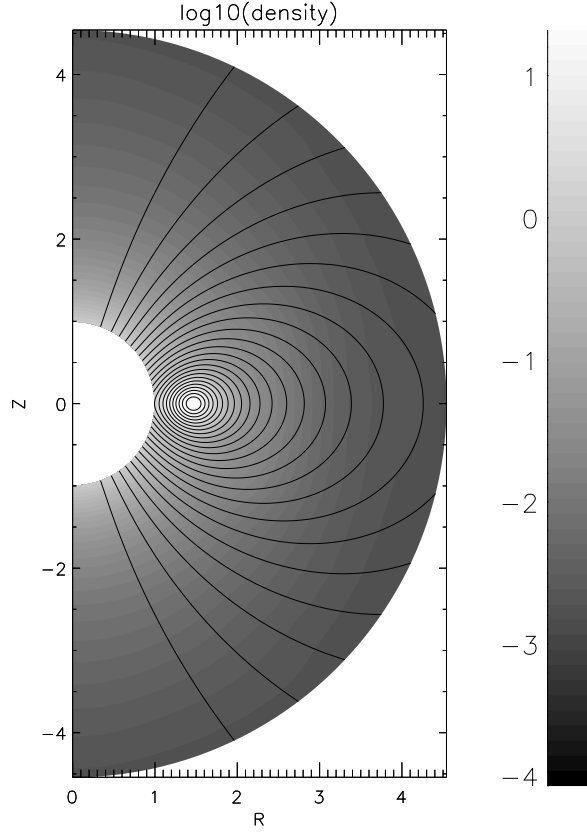


Figure 1. The plasma density $\log_{10} \rho$ (grey scale) and magnetic field lines for the initial relaxed configuration (i.e., at $t = t_0 = 32.0$ or 39.3 hours).

where all quantities except the temperature T are fixed outside the flux emergence region, while T is determined by equivalent extrapolation. The boundary conditions at the poles for the φ - and θ -components of vector quantities are asymmetric, while for the r -components and the plasma density symmetric BCs are applied.

The simulations are done in two phases. In an initial phase the initial state is relaxed for a long period, viz., until $t = 32$ corresponding to 39.3 hours, to assure that a numerical equilibrium is reached with a stable flux rope, see Fig. 1. After this initial phase the kinetic energy of the system and the changes in density and magnetic fields are negligible. Note, that in equations (2) and (3), the current I has been determined by trial and error to make sure that the flux rope center approximately keeps stable for a long enough time, viz., 32 time units. In these simulations the (dimensionless) value of I we obtained this way is 6.33.

In analogy with Forbes *et al.* (1984) and Chen & Shibata (2000), the magnetic flux emergence is then implemented by changing the boundary magnetic field in time. This is done by changing the value of vector potential A_φ starting at $t = t_0 = 32$. The magnetic flux emergence region is limited to $\pi/2 - 0.6 \leq \theta \leq \pi/2 + 0.6$ and grows

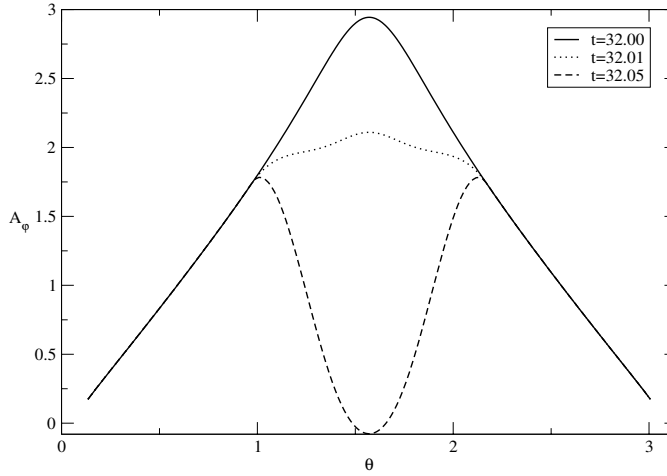


Figure 2. $A_\phi|_{r=1}$ versus θ at $t = t_0 = 32.0$ (after the relaxation), $t = 32.01$ (during the flux emergence), and $t = 32.05$ (after the flux emergence).

linearly in time, in the time interval $\Delta t = t_e - t_0$:

$$A_\phi(t; \theta)|_{r=1} = A_\phi(0; \theta)|_{r=1} + c_e \frac{(t - t_0) \cos^2\left(\left(\frac{\pi}{2} - \theta\right)\frac{\pi}{1.2}\right)}{\Delta t \sin \theta}, \quad t_0 \leq t \leq t_e, \quad (5)$$

where c_e is a free parameter. This BC is illustrated in Fig. 2. After $t = t_e$ all quantities at the bottom boundary are again fixed (except for the value of T) as before.

3. Numerical results

After the initial relaxation of the equilibrium (as described above) the triggering phase starts. In this phase we emerge the magnetic flux at the lower boundary ($r = 1$) in a limited time interval ($\Delta t = t_e - t_0$) and a limited θ -interval around the equator. We considered a range of different time intervals, i.e., different values of the parameter t_e in equation (5). The parameter c_e in equation (5) is set to -3 in all simulations so that the total amount of emerged flux is the same in all simulations, viz., $2\pi c_e \psi_0 \approx -6.6 \times 10^{20}$ Mx in northern hemisphere and the exact opposite in the southern hemisphere. Hence, we effectively change the *flux emergence rate*, i.e., $2\pi c_e \psi_0 / \Delta t$, in this parametric study as the same amount of flux is emerged in a time interval ranging from 0.05 to 3, i.e., from 221 s to 3 h 41 min. The flux emergence rate thus varies from -3×10^{18} Mx/s to -5×10^{16} Mx/s. Note that the emerging magnetic flux has the opposite polarity from the ambient coronal magnetic field.

As the new magnetic flux emerges near the neutral line magnetic reconnections take place below the flux rope leading to partial magnetic cancellation and the loss of equilibrium. Consequently, the flux rope starts moving upwards and the region below the flux rope is evacuated and suffers a pressure drop. Therefore, magnetized plasma at both sides (left and right to the null point) is seen to move inward. A clearly formed current sheet below the flux rope can be seen in Fig. 3. The current in the sheet has the same direction as that in the flux rope and, therefore, flux rope feels a downward Lorentz force because of this sheet current. With a finite resistivity, the current sheet

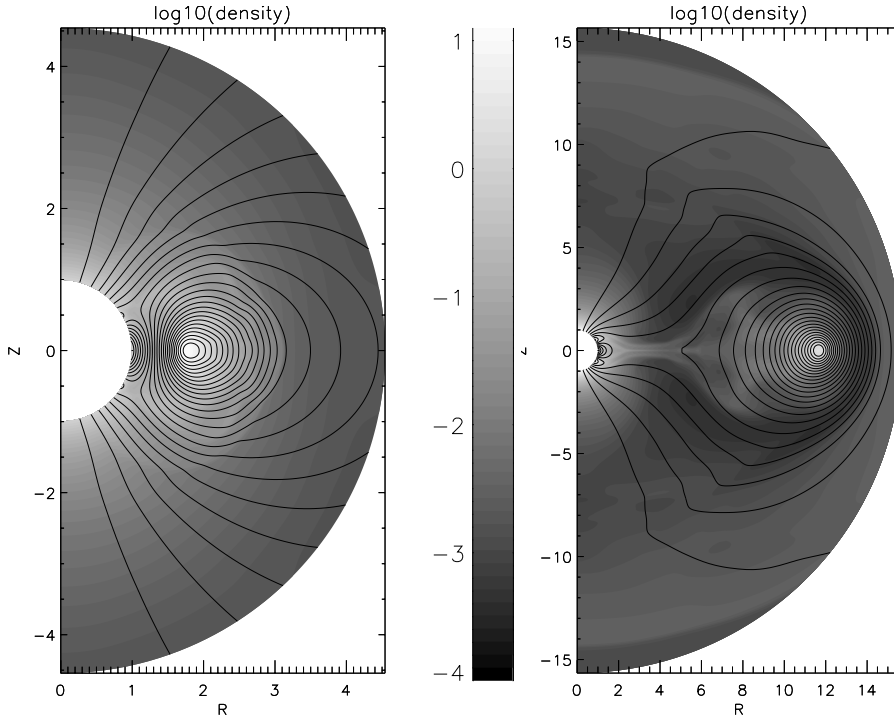


Figure 3. The plasma density $\log_{10} \rho$ (grey scale) and magnetic field lines at $t = 32.3$ (left) and at $t = 36.0$ (right, with a different scale on the abscissa) for the case with $t_e = 32.05$.

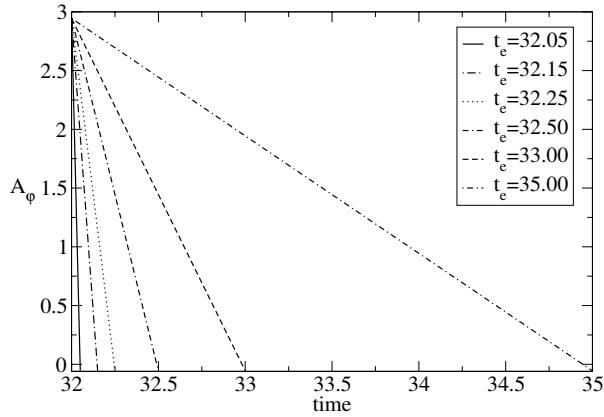


Figure 4. Considered time dependencies of $A_\varphi|_{r=1, \theta=\frac{\pi}{2}}$ for the same c_e value ($= -3$).

would be dissipated and the escaping flux rope would be accelerated in that case. Here however, we only rely on numerical diffusion.

We have performed a parametric study of the emerging flux rate by changing the value of t_e while keeping c_e constant. This is illustrated in Fig. 4. The corresponding

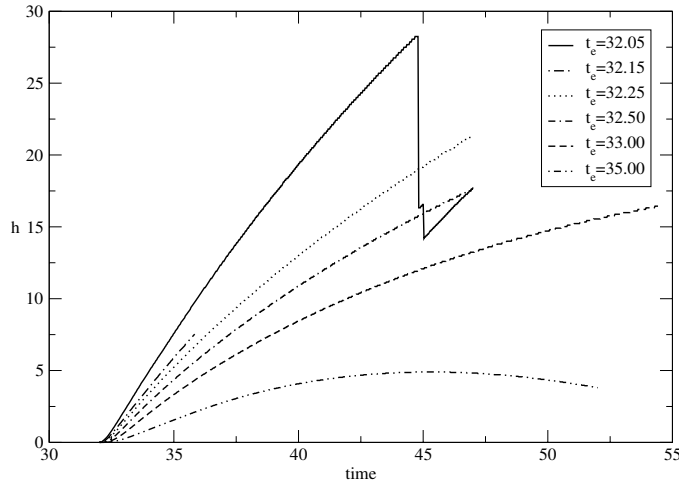


Figure 5. Height of the flux rope center vs. time for the different flux emergence rates.

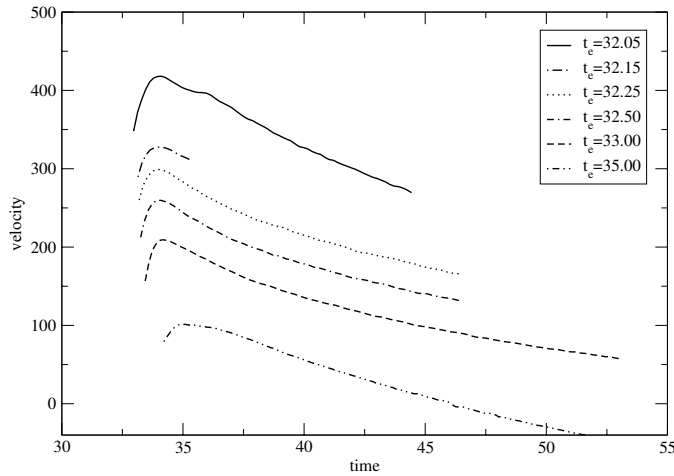


Figure 6. Velocity of the flux rope center vs. time for different flux emergence rates.

height-time curves of flux rope center are shown in Fig. 5. From these curves, the velocities can be easily derived. The related velocities are shown in Fig. 6. Clearly, an increasing flux rate leads to a higher upward speed of the flux rope while the total amount of emerged flux is the same in all cases. The obtained velocities are in the range which is typical for CMEs. For an emerging flux rate that is smaller than some threshold value, the flux rope cannot escape from the corona and it falls down again as a result of the deceleration due to gravity (see e.g., the lowest curve in Fig. 5).

4. Discussion and conclusion

The CME initiation model of Chen & Shibata (2000) has been extended by including the effects of gravity, geometry (curvature), and an ambient medium. Moreover, we studied the effect of the flux emergence rate on the velocity of the resulting flux rope CMEs. The obtained velocities lie in the typical range of 300–400 km/s for reasonable values of the flux emergence rate. For smaller values of the flux emergence rate, the launched flux rope falls back to the Sun as a result of the gravitational deceleration. The present model cannot explain fast CMEs. Yet, these fast CMEs are the important ones for space weather, since these CMEs drive shocks which accelerate particles and impinge on the Earth's magnetosphere. In future analyses we will extend the parametric study to investigate whether we can simulate the deceleration of the fast CMEs and the acceleration of slow CMEs as observed by Gopalswamy (2004). Furthermore, we will include the effect of the drag of the background solar wind.

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Flare Ribbon Expansion and Energy Release

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Abstract. We report a detailed examination about the relationship between the evolution of the H α flare ribbons and the released magnetic energy during the April 10 2001 flare. In the H α images, several bright kernels are observed in the flare ribbons. We identified the conjugated footpoints, by analyzing the lightcurves at each H α kernels, and showed their connectivities during the flare. Then, based on the magnetic reconnection model, we calculated quantitatively the released energy by using the photospheric magnetic field strengths and separation speeds of the H α flare ribbons. Finally, we examined the downward motions which are observed at the H α kernels. We found that the stronger the red-asymmetry tends to be associated with the brighter the H α kernel.

Key words. Sun: activity—Sun: flares, chromosphere, corona, X-rays, gamma rays.

1. Introduction

Based on the CSHKP magnetic reconnection model (Carmichael 1964; Sturrock 1966; Hirayama 1974; Kopp & Pneuman 1976), we can explain several well-known features associated with solar flares (see Fig. 1). The H α flare ribbons, for example, are caused by the precipitation of nonthermal particles and/or the effect of thermal conduction which are produced after magnetic field lines reconnect in the corona. At successive reconnections, the reconnection points (X-points) move upward, and therefore, newly reconnected field lines have their footpoints further out than the footpoints of the field lines that have already reconnected, which leads us to recognize the ‘apparent’ separation motion of the flare ribbons. The H α flare ribbons are located on either side of the magnetic neutral line, and have magnetic polarities opposite to each other. The evolution of the flare ribbons, therefore, has some information on the energy release mechanism.

In this paper we report a detailed examination of the flare ribbon evolution during an X2.3 solar flare which occurred on April 10, 2001. We examined fine structures inside

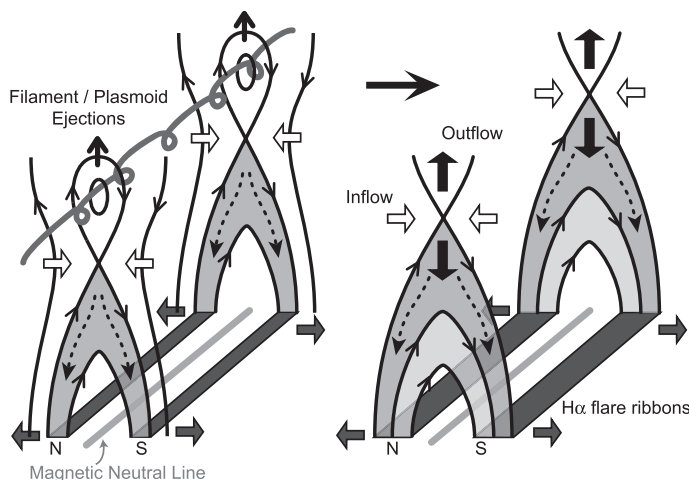


Figure 1. Cartoon of magnetic reconnection. Solid lines show the magnetic field lines. N and S show the magnetic polarities. Reconnected magnetic field lines form post-flare loops with a cusp shaped structure (gray regions). Along the reconnected loops, nonthermal particles and thermal conduction propagate from the reconnection site to the footpoints (dashed arrows). As they precipitate into the chromosphere the $H\alpha$ flare ribbons are generated at the footpoints (dark gray region). They are located on either side of the magnetic neutral line (light gray line), and have magnetic polarities opposite to each other.

flare ribbons by using the $H\alpha$ images. We identified the conjugate footpoints of each $H\alpha$ kernel by using cross-correlation functions of the light curves, and examined their temporal evolution. Next, we estimated the amount of the released magnetic energy quantitatively by using the separation speed of the flare ribbons and the photospheric magnetic field strength. Then, we examined the downward motions observed at the $H\alpha$ kernels. The precipitation pushes the chromospheric plasma downward and modulates $H\alpha$ emission line at the footpoints of the reconnected loops, which is observed as red-shifted emission line, called a “red-asymmetry”. We investigated when and where strong red-asymmetry features.

2. Observations

We observed a typical two-ribbon flare (X2.3 on the GOES scale) which occurred in the NOAA Active Region 9415 (S22°, W01°) at 05:10 UT, April 10, 2001 with the 18 cm Sartorius Refractor (*Sartorius*) at Kwasan Observatory, Kyoto University. The $H\alpha$ monochromatic images of the flare were obtained with the temporal and spatial resolution of about 1 s and 1'', respectively. During the observation, the exposure time was properly regulated so that the fine structure inside the flare ribbons, like $H\alpha$ kernels, can be clearly seen without saturating. The flare was also observed with 60 cm Domeless Solar Telescope (DST) at Hida Observatory, Kyoto University. $H\alpha \pm 0.0 \text{ \AA}$, $+1.5 \text{ \AA}$, and $-1.5 \text{ \AA}$ images were taken with the temporal and the spatial resolutions of about 30 s and about 0.6'', respectively. The extreme-ultraviolet (EUV) images of the flare were obtained with *Transition Region and Coronal Explorer* (TRACE; Handy *et al.* 1999; Schrijver *et al.* 1999). TRACE 171 Å images clearly show post-flare loops which

confine ~ 1 MK-plasma. The flare ribbons in the *TRACE* 171 Å images nearly simultaneously appear with the $H\alpha$ flare ribbons, while the formation of the post-flare loops are noted after the evolution of the flare ribbons. This is because it takes a certain time for hot plasma, which is heated up to about 10 MK or more, to cool down to 1 MK.

To examine the locations of the HXR radiation sources where strong energy releases occur, and to compare the temporal variations of the estimated energy release rates with the light curve of the HXR total intensity, we used HXR data taken with the hard X-ray telescope (HXT; Kosugi *et al.* 1991) aboard *Yohkoh* (Ogawara *et al.* 1991). The temporal resolution of the HXT data is 0.5 s. We also used microwave time profiles obtained with the Nobeyama Radioheliograph (NoRH; Nakajima *et al.* 1994). The NoRH microwave data were obtained at 17 GHz with a temporal resolution of 1.0 s. Both the HXT and NoRH light curves show some bursts. The dawn of *Yohkoh* was at 00:19 UT, and the earlier part of the flare was missed because of night. No large NoRH bursts are seen from 00:20 to 00:31 UT, since the emission mechanism becomes optically-thick during the time range, while the other bursts are from optically-thin gyrosynchrotron radiation. Photospheric magnetograms were taken with the Michelson Doppler Imager (MDI; Scherrer *et al.* 1995) onboard the *Solar and Heliospheric Observatory* (SOHO; Domingo *et al.* 1995). We co-aligned the $H\alpha$ images and the MDI images visually. The accuracy of the co-alignment is about $2''$, which is comparable to the pixel size of the MDI data.

3. Identification of conjugate footpoints

Firstly, we identified the conjugate pairs of the footpoints seen in the $H\alpha$ images by a new method which Asai *et al.* (2003) suggested. Figure 2(a) and (b) show a photospheric magnetogram taken with the MDI and an $H\alpha$ image obtained with *Sartorius*, respectively. We divided the $H\alpha$ flare ribbons, which have the magnetic polarities opposite to each other, into fine meshes (see Fig. 2c). Then, we drew light curves of total intensity for each box in both the meshes, and identified the highly-correlated pairs (conjugated pairs) by calculating the cross-correlation functions. In Fig. 2(c) *black solid* lines are drawn to connect such highly-correlated pairs.

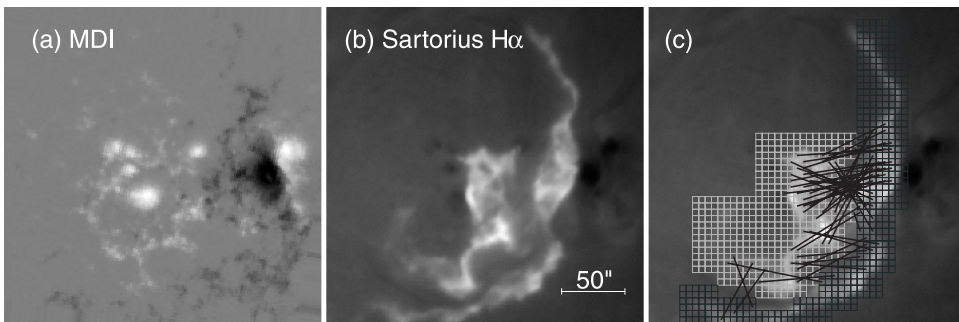


Figure 2. Method of analyses. Celestial north is up, and west is to the right. (a) Photospheric magnetogram taken with *SOHO* MDI. (b) Raw $H\alpha$ image of the *Sartorius* at 05:31 UT. (c) $H\alpha$ image overlaid with the fine meshes and with black solid lines which connect highly-correlated pairs on (b). The left (east) mesh lines is positive polarity (drawn with light gray lines), and the right (west) one (dark gray lines) is negative polarity.

Then, we examined the spatial relationship between the post-flare loops seen in the EUV images and the conjugated pairs of the $H\alpha$ kernels. We checked whether the $H\alpha$ pairs are actually connected by the $171\text{ \AA}$ flare loops or not, and found that almost all the pairs are connected by the flare loops. This good coincidence between the $H\alpha$ pairs and the flare loops strongly supports the scenario of the flare loop formation: the EUV flare loops are formed by the plasma evaporated from the brightening $H\alpha$ pairs. We also examined, in detail, the temporal and spatial evolution of the pairs of the $H\alpha$ conjugate footpoints selected above. The start times of brightening of each $H\alpha$ pair tell us the site and the time of the energy release or the magnetic field reconnection. Thus, we follow the evolution of the energy release sites by examining the fine structures inside the flare ribbons or $H\alpha$ kernels. At the beginning, the $H\alpha$ kernels and the conjugated pairs appear nearly parallel to the magnetic neutral line or are in a strongly sheared configuration. As the flare progresses, the flare ribbons quickly separate from each other, and many $H\alpha$ kernels appear successively. The locations of the $H\alpha$ kernels tend to move toward northeast in the eastern ribbon and southwest in the western ribbon. Such changes from the strongly-sheared to less-sheared directions of the lines that are connecting $H\alpha$ pairs indicate the configuration of the height-dependent sheared magnetic field in the corona over the neutral line.

4. Energy release rate

Secondly, we estimated the amount of energy released via magnetic reconnection during the flare (see Asai *et al.* (2004) for more details). Based on the reconnection models, the energy release rate can be written as the product of the Poynting flux into the reconnection region S ($= 2(4\pi)^{-1}B_c^2v_i$) and the area of the reconnection region A as follows:

$$\frac{dE}{dt} = SA = 2\frac{B_c^2}{4\pi}v_iA, \quad (1)$$

where B_c is the magnetic field strength in the corona and v_i is the inflow velocity into the reconnection region. Note that this assumption is based on the 2D expression of the magnetic reconnection model. It is very difficult to measure the coronal magnetic field strength B_c and the inflow velocity v_i directly. Consequently, an indirect method is needed to evaluate the energy release rate quantitatively by using observable values. We assume that the area of the reconnection region A does not change significantly during the flare and is independent of the magnetic field strength, and that the energy release rate is simply proportional to the Poynting flux S .

Here, we examined the relation between the expanding motions of the $H\alpha$ flare ribbons and the released magnetic energy by using the magnetic field strengths at the photospheric level B_p and the separation speed of the flare ribbons v_f , instead of direct coronal values such as B_c and v_i . We estimated the reconnection rate Φ , i.e., the electric field strength in the corona. The reconnection rate is one of the most important physical values in reconnection physics, because it shows how violently magnetic reconnection progresses. It is defined as the reconnected magnetic flux per unit time, and expressed as $\Phi = B_c v_i$. From the conservation of magnetic flux, we can write $\Phi = B_p v_f$. We also estimated the Poynting flux S . If we assume that the area of the reconnection region (A) does not change significantly during the flare and is independent of the

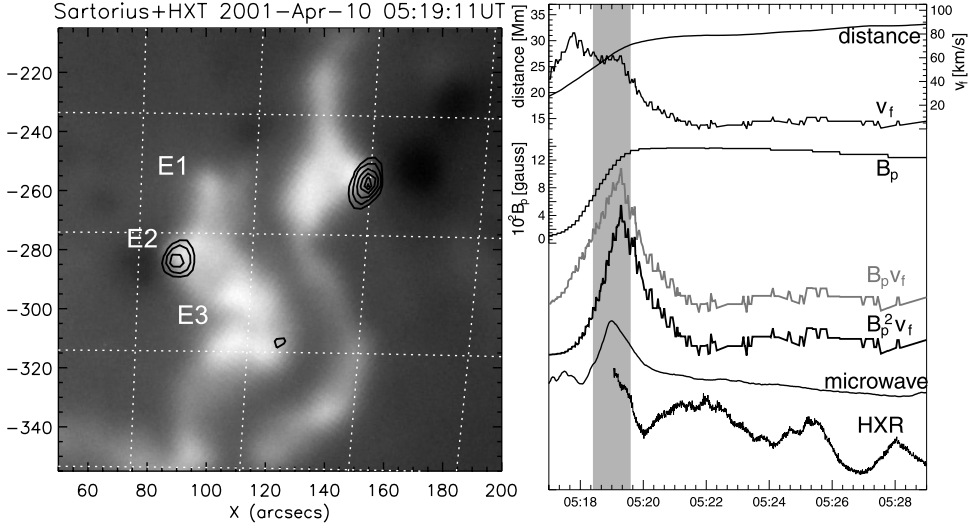


Figure 3. **Left:** $H\alpha$ image kernels overlaid with HXR contour image. E1, E2, and E3 show $H\alpha$ kernels. Only E2 is associated with the HXR emission source. **Right:** Temporal variations of the physical parameters measured along the slit line which passes through the E2 $H\alpha$ kernel. **From top to bottom:** distance of the outer edge of the ribbon from the magnetic neutral line; separation speed of the ribbon (v_f); photospheric magnetic field strength (B_p) at the outer edge of the ribbon; estimated reconnection rate Φ (gray lines); estimated Poynting flux S ; microwave correlation plot in 17 GHz obtained with NoRH; HXR count rate measured in the Yohkoh/HXT H band (53–93 keV).

magnetic field strength, the S is written as $2(4\pi)^{-1}B_c^2v_i$. Furthermore, it is rewritten as $S \propto B_p^2v_f$ with the assumption that B_c is proportional to B_p in the same ratio all over the flaring region. Although the assumption may be simplistic, it is reasonable, provided we do not discuss the energy release rates of the whole active region. This is because the B_c -to- B_p ratio remains approximately constant in narrow regions.

In the April 10, 2001 flare, the expansion motion of the flare ribbons is roughly seen in the east–west direction, since the magnetic neutral line lies approximately in the celestial north–south direction (see Fig. 2b). Therefore, we put slit lines in the east–west direction so that each passes the flare ribbons perpendicularly, and tracks the positions of the outer edges of the ribbons along the slit lines. In the right panel of Fig. 3, we show the temporal variation of the physical parameters, such as B_p , v_f , Φ , and S along a slit line. Here, we defined v_f as the speed of the flare-ribbon separation, relative to the magnetic neutral line, in the perpendicular direction. The right panel of Fig. 3 also shows the estimated Φ and S compared with the light curves in HXR and microwaves which were obtained with HXT and NoRH. The evolutions of the reconnection rates and Poynting fluxes appear to match the HXR and microwave light curves well (see vertical light gray lines).

Are the estimated energy release rates at the $H\alpha$ kernels associated with the HXR bursts large enough to explain the difference between the spatial distribution of the HXR sources and that of the $H\alpha$ kernels? To answer the question, we compared them with those at the $H\alpha$ kernels without HXR bursts, e.g., E1 and E3 in the left panel of Fig. 3, at the time of the first HXR burst (05:19UT). We found that the reconnection

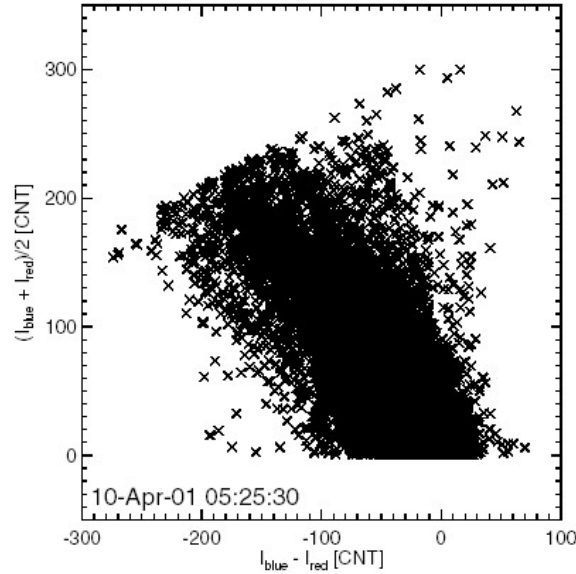


Figure 4. Scatter plot between the red-asymmetries and the intensities of the $H\alpha$ kernels. The horizontal axis shows the differential between I_{blue} and I_{red} . The vertical axis shows their averaged intensity.

rate of E2 is at least 16 times stronger than that of the other two kernels. The Poynting flux at E2 is also much more enhanced, and is 150 times stronger than that of the others. Both the estimated reconnection rates and Poynting fluxes are large enough locally at the HXR sources. The results can explain the difference between the spatial distribution of the HXR sources and that of the $H\alpha$ kernels.

5. Red asymmetry

Finally, we examined the evolution of the $H\alpha$ kernels spectroscopically. At $H\alpha$ kernels, the precipitation of the nonthermal particles and/or thermal conduction into the chromosphere generate the downward compression of the chromospheric plasma, which is observed as the reddening of $H\alpha$ emission (Ichimoto & Kurokawa 1984). The authors also reported that the red-shift corresponds to the downward motion with a speed of about $50\text{--}100\text{ km s}^{-1}$, based on their spectroscopic observations.

We measured the differential between the $H\alpha$ intensity in the red-wing ($H\alpha + 1.5\text{ \AA}$) and that in the blue-wing ($H\alpha - 1.5\text{ \AA}$), that is, how bright are the red-wing images by using the $H\alpha$ wing data obtained with DST, and examined the spatial distribution of the red-asymmetry. We found that the red-asymmetry appears all over the flare ribbons during the impulsive phase of the flare, and above all it is conspicuous at the outer edges of the flare ribbons. We also examined the relationship between the red-asymmetry and the intensity of the $H\alpha$ kernels. Figure 4 shows the scatter plot between the red-asymmetry ($I_{\text{blue}} - I_{\text{red}}$; horizontal axis) and the averaged $H\alpha$ intensity $((I_{\text{blue}} + I_{\text{red}})/2$; vertical axis) at the kernels. We can see the tendency that the stronger red-asymmetry (brighter in the red-wing images) is associated with the brighter $H\alpha$ kernel.

6. Summary and discussions

We examined in detail the fine structure and the temporal and spatial evolution of the H α flare ribbons associated with the April 10, 2001 flare.

We identified the conjugate footpoints in both the flare ribbons by calculating cross-correlation functions, and confirmed that the pairs of H α conjugate footpoints are really connected by the post-flare loops seen in the EUV images. This result strongly supports the flare scenario that the H α kernels are brightened by the precipitation of high energy electrons or by heat conduction and that evaporated plasma from the H α kernels produce the high-temperature flare loops observed in SXR and EUV. Our finding suggests that we can study the magnetic field configuration in the energy release sites by examining evolution of H α kernels. We have been able to follow the formation of the flare loops and the evolution of energy release sites. Next, we estimated the energy release rate quantitatively, based on the magnetic reconnection model, by using the photospheric magnetic field strength and the separation speed of the H α flare ribbons. The estimated energy release rates of the H α kernels associated with the HXR sources are locally large enough to explain the difference between the spatial distribution of the H α kernels and the HXR sources. The temporal evolution of the energy release rates also shows peaks corresponding to the HXR bursts. Then, we examined spatially resolved red-asymmetry distributions and their temporal evolution, by using H α wing images. Red-asymmetry tells us the quantitative information of the precipitation (nonthermal particles/thermal conduction). We found the tendency that the stronger red-asymmetry is associated with the brighter H α kernel.

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Solar X-ray Spectrometer (SOXS) Mission – Low Energy Payload – First Results

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Abstract. We present the first results from the ‘Low Energy Detector’ payload of ‘Solar X-ray Spectrometer (SOXS)’ mission, which was launched onboard GSAT-2 Indian spacecraft on 08 May 2003 by GSLV-D2 rocket to study the solar flares. The SOXS Low Energy Detector (SLD) payload was designed, developed and fabricated by Physical Research Laboratory (PRL) in collaboration with Space Application Centre (SAC), Ahmedabad and ISRO Satellite Centre (ISAC), Bangalore of the Indian Space Research Organization (ISRO). The SLD payload employs the state-of-the-art solid state detectors viz., Si PIN and Cadmium-Zinc-Telluride (CZT) devices that operate at near room temperature (-20°C). The dynamic energy range of Si PIN and CZT detectors are 4–25 keV and 4–56 keV respectively. The Si PIN provides sub-keV energy resolution while CZT reveals ~ 1.7 keV energy resolution throughout the dynamic range. The high sensitivity and sub-keV energy resolution of Si PIN detector allows the measuring of the intensity, peak energy and equivalent width of the Fe-line complex at approximately 6.7 keV as a function of time in all 8 M-class flares studied in this investigation. The peak energy (E_p) of Fe-line feature varies between 6.4 and 6.8 keV with increase in temperature from 9 to 34 MK. We found that the equivalent width (w) of Fe-line feature increases exponentially with temperature up to 20 MK but later it increases very slowly up to 28 MK and then it remains uniform around 1.55 keV up to 34 MK. We compare our measurements of w with calculations made earlier by various investigators and propose that these measurements may improve theoretical models. We interpret the variation of both E_p and w with temperature as the changes in the ionization and recombination conditions in the plasma during the flare interval and as a consequence the contribution from different ionic emission lines also varies.

Key words. Solar Flares—X-ray detectors—X-ray line emission—equivalent width.

1. Introduction

The ‘Solar X-ray Spectrometer (SOXS)’ mission (Jain *et al.* 2000a, b; 2005a, b) was launched onboard an Indian geostationary satellite namely GSAT-2 on 08 May 2003 by GSLV-D2 rocket. The SOXS aims to study the high energy and temporal resolution X-ray spectra from solar flares. The SOXS consists of two independent payloads, viz.,

SOXS Low Energy Detector (SLD) and SOXS High Energy Detector (SHD). The SLD is comprised of two semiconductor devices, viz., Silicon PIN detector for 4–25 keV (area 11.56 mm²); and Cadmium Zinc Telluride (CZT) detector for 4–56 keV energy range (area 25 mm²). These state-of-the-art solid state detectors in SLD operate at near room temperature, i.e., at -20°C . Both detectors have 100 ms temporal resolution characteristics, which make them most appropriate for solar flare research in the context of energy transport and acceleration time scales of particles. The energy resolution revealed by Si PIN detector is sub-keV while CZT detector, however, could provide 1.7 keV throughout their dynamic energy range. The SLD payload is designed and developed at the Physical Research Laboratory (PRL) in collaboration with ISRO Satellite Centre (ISAC), Bangalore, and Space Application Centre (SAC), Ahmedabad.

It has been widely established that the solar corona below 5 keV is so hot that a large number of photons are emitted, mostly in the form of lines, even in the absence of the flare. This restricts the design of the spectroscopy experiment from 1 keV to 20 keV in view of saturation of the detector, because of its limited counts handling capability, by low energy photons only. However, in order to improve our current understanding on the X-ray line emission characteristics the synoptic observations at energies below 10 keV are of utmost importance, which may reveal the temperature enhancement during flares of different magnitude. On the other hand, it has been shown by Jain *et al.* (2000a, b, 2005a, b) that iron complex lines (Fe xxv, xxvi) at 6.7 keV and Fe/Ni complex lines at 8 keV appear only during solar flare activity and understanding of their emission characteristics requires extremely high spectral and temporal resolution observations. This requires design of high spectral and temporal resolution X-ray spectrometer but in order to avoid saturation of the detectors it needs optimization between resolution and satisfactory performance of the detectors during all classes of flares.

The simulation of X-ray emission from the solar flare (Jain *et al.* 2000a, b, 2005a) unambiguously indicates the possibility to determine the thermal spectrum from the measurements of the soft X-ray line emission. Thus measurements of soft X-ray flux before and during the flare provide a wonderful opportunity to study the soft X-ray characteristics of active region corona. The high sensitivity and sub-keV energy resolution of Si PIN detector allows the intensity and mean energy of the Fe-line complex at approximately 6.7 keV to be measured as a function of time in all classes of flares. This line complex is mostly due to the 1 s-2 p transitions in He-like and H-like iron, Fe xxv and Fe xxvi respectively, with associated satellite lines. Another weaker line complex at ~ 8 keV made up of emission from He-like nickel and more highly excited Fe xxv ions is also evident in the more intense flares (Phillips 2004; Phillips *et al.* 2004). Detailed calculations of emission line intensities as a function of temperature, with provision for different element abundance sets (e.g., photospheric or coronal), are given by the MEKAL/SPEX atomic codes (Mewe *et al.* 1985a, b; Phillips *et al.* 2004) and the CHIANTI code (Dere *et al.* 1997). These codes also include thermal continuum intensities. These codes are used to interpret the SLD spectral observations in terms of the plasma temperature and emission measure. The centroid energy and width of the iron-line complex at ~ 6.7 keV, the intensity of the Fe/Ni line complex at ~ 8 keV, and the line-to-continuum ratio are the functions of the plasma temperature and can be used to limit the range of possible plasma parameters. However, detailed study of such features of the Fe and Fe/Ni line complexes has not been carried out mainly due to non-availability of spectral observations in the energy range

3–10 keV and in particular with high spectral and temporal resolution, which are critically required to measure precisely the line features and plasma parameters. The high spectral and temporal resolution spectra may reveal many unidentified lines as shown by RESIK Bragg crystal spectrometer aboard CORONAS-F (Sylwester *et al.* 2004). Phillips *et al.* (2004) carried out a study of solar flare thermal spectrum using RHESSI, RESIK and GOES mission data and determined absolute elemental abundances, which however may have subjected to uncertainties due to measurements from three different instruments that were not calibrated by a single common technique. However, the SOXS mission provides the X-ray spectra in the desired 4–10 keV energy band with improved spectral and temporal resolution. Therefore the purpose of this paper is to study the X-ray emission characteristics of Fe-line feature in solar flares using the high sensitivity and sub-keV energy resolution capabilities of Si PIN detector of SOXS mission. We present the current study of the Fe-line emission as the first results from the observations made by the SLD/SOXS mission.

2. Observations

2.1 Brief instrumentation

The instrumentation of the SLD payload, its in-flight calibration and operation has been described by Jain *et al.* (2005a). However, a brief description of the experiment is as follows: The “Solar X-ray Spectrometer (SOXS)” mission (Jain *et al.* 2000a, b, 2003, 2005a, b) was launched onboard an Indian geostationary satellite namely GSAT-2 on 08 May 2003 by GSLV-D2 rocket. The SLD/SOXS aim to study the high energy and temporal resolution X-ray spectra from solar flares. The SLD payload is functioning satisfactorily onboard the GSAT-2 spacecraft and so far more than 300 flares of importance than GOES C1.0 have been observed. The SLD is comprised of two semiconductor devices, viz., Silicon PIN detector for 4–25 keV (area 11.56 mm²); and Cadmium Zinc Telluride (CZT) detector for 4–56 keV energy range (area 25 mm²). The spectral resolution revealed by Si detector is 0.7 keV @ 6 keV and 0.8 keV @ 22.2 keV, which is better than the earlier detectors used for solar flare research in this energy range. However, spectral resolution achieved from CZT detector is poor, i.e., almost 1.7 keV but it remains stable throughout its dynamic energy range of 4–56 keV. Further, their temporal resolution capabilities are also superb, however we designed for 100 ms during flare mode in order to achieve feasible energy spectrum. The critical operating temperature of both the detectors in the range –5 to –30°C is achieved using thermoelectric cooler that coupled with the detector. The detector package is mounted on a Sun Aspect System which keeps the Sun in the center of the detector for an interval between 03:40 and 06:40 UT everyday. However, after 06:40 UT the temperature on the detectors exceeds the limit to cool down by thermoelectric cooler. Thus SLD payload provides uninterrupted data for ~3 h everyday. The SLD data are of two types – temporal mode (light curves) and spectral mode. The in-flight tests and onboard calibration of the detectors were successfully carried out between 22 May and 06 June 2003. The onboard calibration is carried out by Cd¹⁰⁹ radioactive source mounted inside the collimator of each detector. Thus by integrating the spectra over an hour may yield a line at 22.2 keV for onboard calibration of the detector. So far no anomaly is observed except loss of 14 channels in CZT detector that was observed

Table 1. SLD/SOXS flare events considered for investigation.

Sl. no.	Date	Time UT			GOES Class	Active region	
		Begin	Peak	End		Location	NOAA
1.	30 July 2003	0407	0409	0428	M2.5	N16 W55	10422
2.	13 November 2003	0454	0501	0510	M1.6	N04 E85	10501
3.	19 November 2003	0358	0402	0419	M1.7	N01 E06	10501
4.	07 January 2004	B0355	0400	0433	M4.5	N02 E82	10537
5.	25 March 2004	0429	0438	0507	M2.3	N12 E82	10582
6.	25 April 2004	0528	0536	0558	M2.2	N13 E38	10599
7.	14 August 2004	0413	0414	0432	M2.4	S13 W30	10656
8.	31 October 2004	0526	0531	0546	M2.3	N13 W34	10691

during the first in-orbit test. Both detectors are operating satisfactorily and data are made available daily to PRL for further processing.

2.2 Data set

The first light from the Sun was fed into the detectors on 08 June 2003. The flare trigger threshold was intentionally kept higher so as to observe the signal in contrast to background. The temporal data, i.e., intensity (counts/s) as a function of time is revealed in four energy band viz., 6–7 keV (L1), 7–10 keV (L2), 10–20 keV (L3) and 4–25 keV (T) by Si detector, while in five energy bands by CZT detector viz., 6–7 keV, 7–10 keV, 10–20 keV, 20–30 keV and 30–56 keV. In Table 1 we show the flare events analyzed by us to study the X-ray spectral evolution of Fe-line feature in the flare plasma. We selected eight flares of *GOES* importance class M for the current study as first results. However, in preview to the goal of studying Fe-line feature, we use data from Si detector only for the current investigation.

2.2.1 Temporal mode

In Fig. 1 we show the temporal mode observations, i.e., light curves of 31 October 2004 flare in four energy windows of Si detector. The time resolution for temporal and spectral mode observations during quiet period is 1 s and 3 s respectively but during flare it is 100 ms for both temporal and spectral modes. The intensity (counts/s) of the light curve shown in Fig. 1 is 20 s moving average of the 100 ms observed data. It may be noted that the flare is gradually rising and long enduring.

2.2.2 Spectral mode

The energy region 4–15 keV in solar flare X-ray spectrum is of great importance in inferring the properties of the hottest parts of the thermal plasma created during a solar flare. It contains emission lines of highly ionized Ca, Fe, and Ni atoms and a continuum that falls off steeply with increasing energy. In this context, SLD is the first payload, which has a dedicated dynamic energy range of 4–25 keV to study the line emission and continuum with sub-keV spectral resolution. This is achieved by employing Si PIN detector as described in the preceding section.

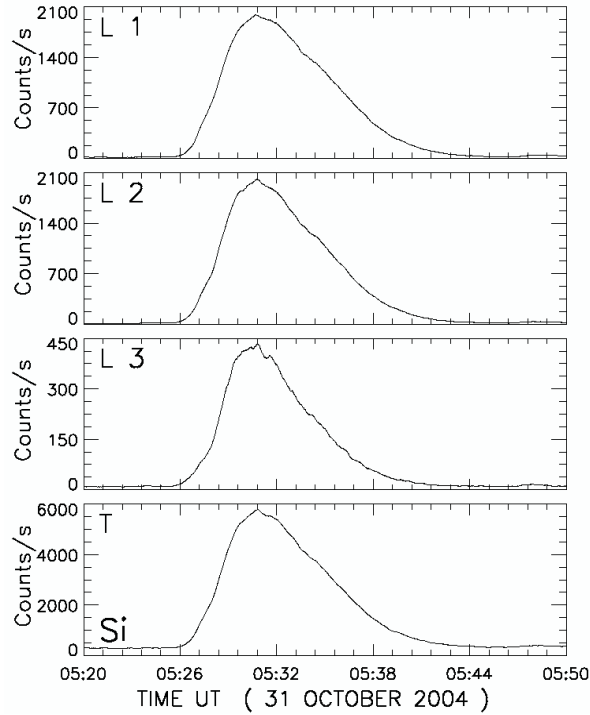


Figure 1. Light curves of 31 October 2004 solar flare as recorded in L1, L2, L3 and T energy bands (see text) of Si detector of SLD/ SOXS mission.

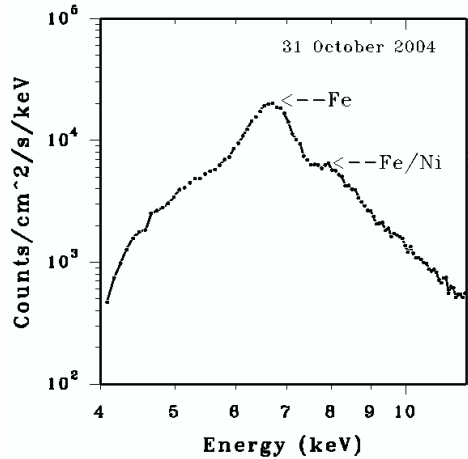


Figure 2. Count spectra from Si PIN detector for 31 October 2004 flare at 05:30:59 UT. Note Fe and Fe/Ni line features.

The energy spectrum, intensity (counts/s) as a function of energy at a given time, in the energy range 4–25 keV distributed over 256 channels with channel width of 0,082 keV, obtained from the instrument is in the form of count spectra, the Si detector's count spectra at the peak time of 31 October 2004 flare is shown in Fig. 2. The low

intensity below 6 keV is due to aluminum plus kapton filter mounted on the detector head to cut the X-ray photons up to 4 keV and electrons up to 300 keV falling in the line-of-sight of the detector (Jain *et al.* 2005). It may be noted that in Si count spectra the Fe and Fe/Ni lines are unambiguously visible at ~ 6.7 and ~ 8 keV respectively. The count spectra are de-convoluted over the instrumental response to obtain the photon spectra, which are indeed useful to study the X-ray line and continuum emission.

3. Analysis and results

The raw data for temporal and spectral mode observations is first corrected for any spurious or false flare as well as for pre-flare background (Jain *et al.* 2005a). The spectrum at a given time is made by integrating the high cadence (100 ms) spectra over an interval of 30 to 100 s period. The photon spectrum is produced by de-convolution of the count spectrum over the instrumental response. The various steps from data acquisition to data analysis are projected in the flow chart shown in Fig. 3. The photon spectra are used to study the evolution of Fe and Fe/Ni lines in a given flare as a function of time. Each photon spectra is formed by integrating the several spectra observed at the cadence of 100 ms or 3 s interval.

3.1 X-ray emission from Fe-line

In order to study the Fe and Fe/Ni line emission, it is rather more important to study their evolution with the flare development, i.e., as a function of temperature because the line emission and its intensity vary with temperature and emission measure (Phillips 2004). Shown in Fig. 4 is a sequence of photon spectra of 31 October 2004 flare in the energy range 5 to 12 keV. The sequence shows evolution of the Fe and Fe/Ni lines as a function of time. It may be noted from this figure that the peak intensity, peak energy and area under the curve of the lines vary over time. In fact, the plasma temperature and hence the emission measure vary over time and these factors mainly control the shape of the line. However, non-thermal contribution also plays a role but in this paper we consider only temperature and emission measure as important parameters.

The Fe-line feature is here defined as the excess above the continuum, as observed by Si spectrometer with spectral resolution (FWHM) ≤ 0.7 keV, in the energy range 5.8–7.5 keV (Phillips 2004). It may be noted from the temporal evolution of this line shown in Fig. 4 that Fe-line features including the peak energy and intensity vary over the flare evolution, which suggests that the abundance and peak energy of the emission line vary as a function of temperature. In this paper we intend to investigate the variation in peak energy of the Fe-line feature by Gaussian fit, which leads us to measure the central peak energy for a given spectra. We analyzed 10 to 27 spectra, for each of the flares under study, depending on the duration of the flare. SOXSoft package (Patel and Jain 2005) is the software package used for data analysis. SOXSoft is specially developed for SOXS mission for data processing and spectra formation.

Once the photon spectra is formed (cf. Fig. 4) we undertake their analysis for deriving plasma parameters such as temperature, emission measure and spectral index using SOXSoft spectra fit program. This program takes its main routine from Solarsoft where Mewe and Chianti codes can be used to derive the plasma parameters. In order to fit the spectra in the energy range between 5 and 15 keV and particularly in the Fe-line feature by isothermal plasma we use Chianti code because thermal continuum from

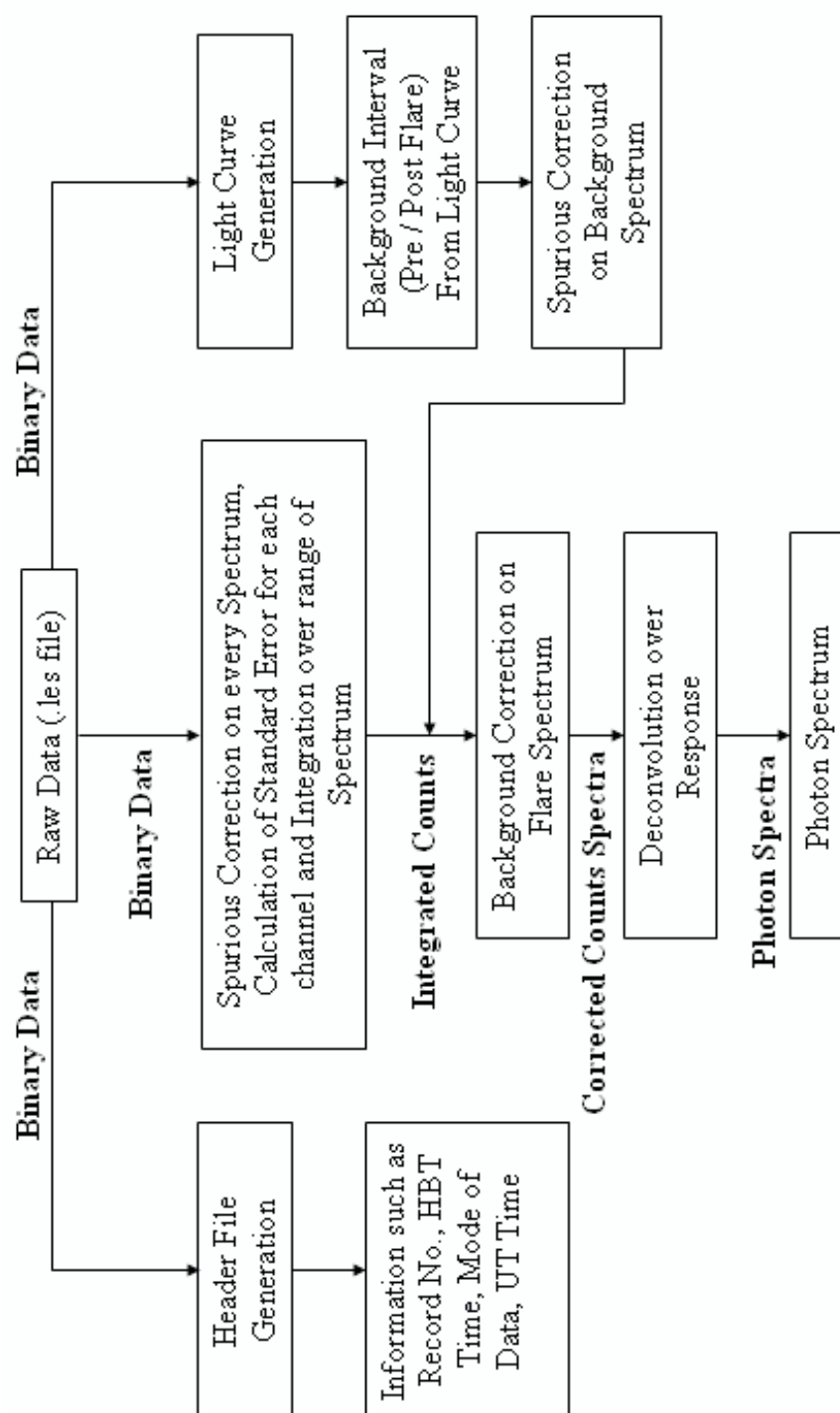


Figure 3. Flow-chart of various steps of SLD/SOXS raw data processing and analysis.

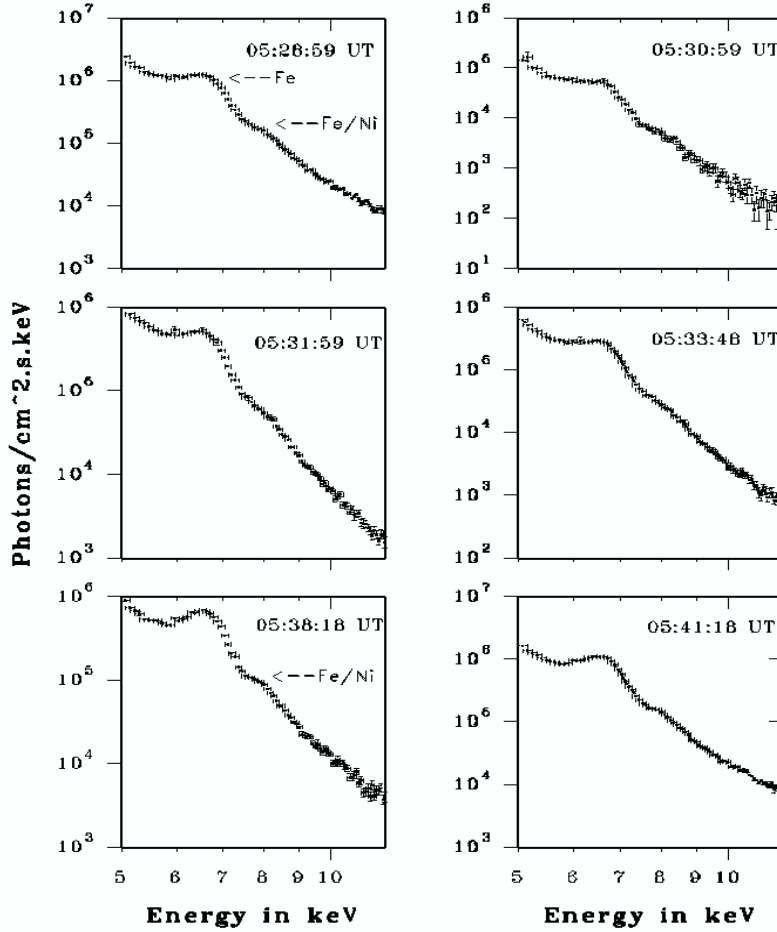


Figure 4. Sequence of X-ray photon spectrum in the energy range 5–12 keV of 31 October 2004 flare showing evolution of Fe and Fe/Ni line features. X-axis error bar is channel width of 0.082 keV, while Y-axis error bar is $\pm 1\sigma$ of the photon flux in the given channel.

it is within 1% of the detailed calculations of Culhane (1969) and the approximation of Mewe *et al.* (1985a). We use the best-fit to the line feature based on the minimum reduced χ^2 (difference counts). In order to derive the line parameters such as the peak energy (E_p), net area and gross area under the curve and equivalent width we subtracted the continuum contribution to the spectrum. The temperatures are derived from the continuum in the energy range 9.5 to 16 keV using Chianti best-fit code. A flow chart of the steps of line analysis and hard X-ray spectra analysis is shown in Fig. 5.

3.1.1 Evolution of temperature and emission measure

Temperatures are derived from the continuum part in the photon spectrum. In Fig. 6 we show a best-fit of photon flux from Chianti code of isothermal plasma temperature and emission measure for the energy range generally between 9.5 and 16 keV for a photon spectra of 31 October 2004. The temperature and emission measures for this

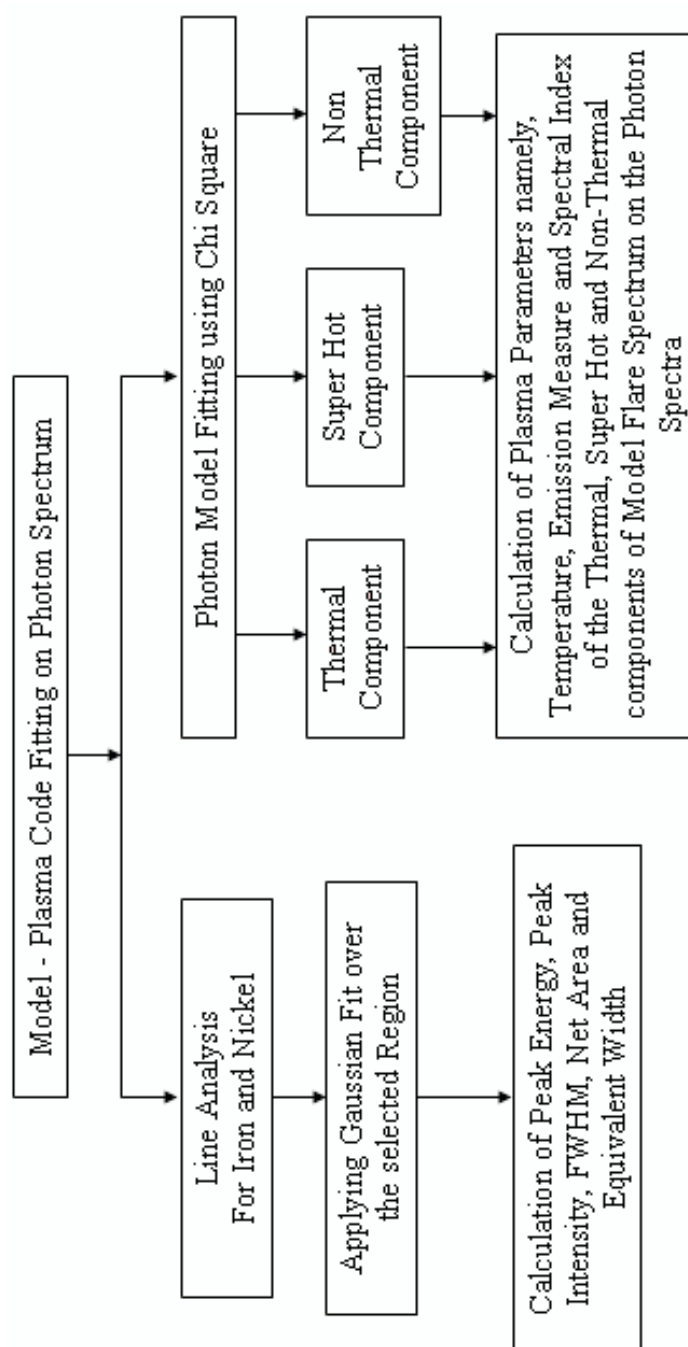


Figure 5. Flow chart of steps of data reduction in order to measure the plasma parameters and Fe-line parameters.

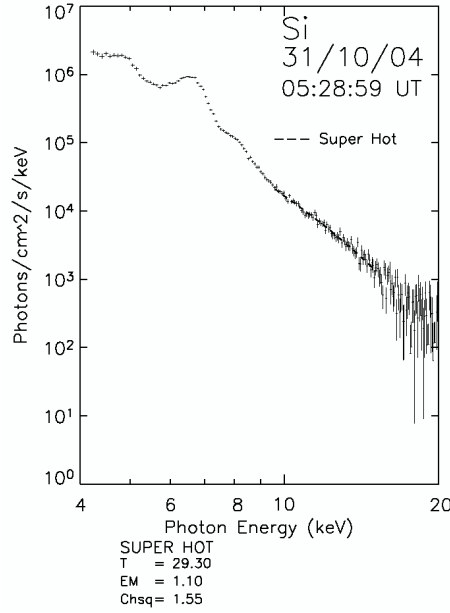


Figure 6. The X-ray photon spectra of 31 October 2004 at 05:28:59 UT. Note the 9.5–16 keV continuum fit by isothermal plasma temperature and emission measure. X-axis error bar is channel width of 0.082 keV, while Y-axis error bar is $\pm 1\sigma$ of the photon flux in the given channel.

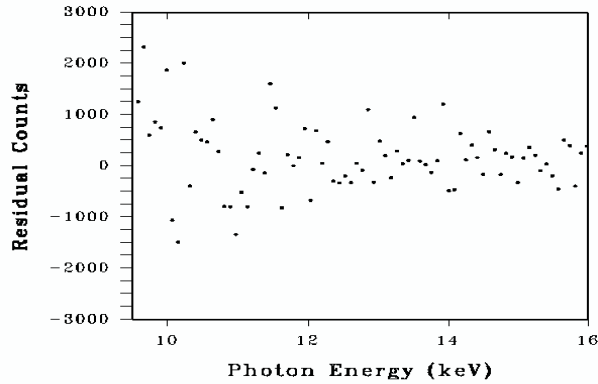


Figure 7. The residual (difference) counts of isothermal plasma continuum fit (cf. Fig. 6).

spectrum are 29.3 MK and $1.3 \times 10^{49} \text{ cm}^{-3}$ respectively. The isothermal fit by Chianti code using Solarsoft is accepted if reduced $\chi^2 < 5$. For example the residual counts for the continuum fit for a χ^2 of 1.55 (Fig. 6) are shown in Fig. 7. In this way, we obtain temperature and emission measure values for each photon spectra of a given time of the flare. We studied almost 10 to 27 photon spectra for each flare depending upon its duration. In Fig. 8, as an example, we show the temperature and emission measure evolution for the 31 October 2004 flare. It may be noted that the evolution of the temperature is almost similar to the light curve of the flare (Fig. 1), and peaking around flare maximum. This may indicate that in this flare X-ray photon emission was

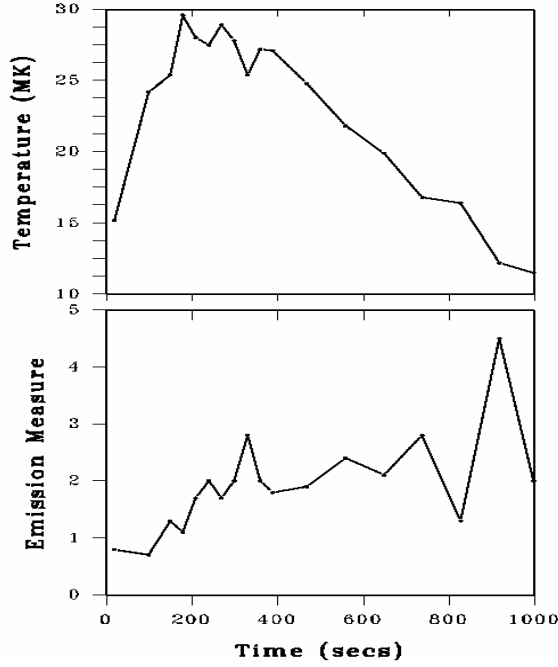


Figure 8. Evolution of temperature and emission measure derived from the continuum of the X-ray photon spectra of 31 October 2004 flare.

more governed by temperature of the plasma. On the other hand, emission measure was between 1 and $3 \times 10^{49} \text{ cm}^{-3}$ throughout the flare except in one spectrum in which it was seen to be higher than $4 \times 10^{49} \text{ cm}^{-3}$.

3.1.2 Peak energy of Fe-line features

The thermal component in Si PIN spectra is observed to have a prominent broadened emission line feature at 6.7 keV and a less intense line feature at 8 keV indicating high plasma temperatures. The 6.7 keV feature corresponds to a group of emission lines due to Fe xxv, associated dielectronic satellites of ions from Fe xix to Fe xxiv, and fluorescence-formed lines of Fe ii, and a second group due to Fe xxvi ($\text{Ly}\alpha$) lines and associated satellites. The Fe xxv lines are excited at electron temperatures $T_e \geq 12 \text{ MK}$, while the Fe xxvi lines are excited at $T_e \geq 30 \text{ MK}$ (Phillips 2004). This line complex is referred to in this paper as the Fe line feature. Thus we may conclude that the Fe-line feature is made up of many individual lines each having its own temperature dependence. Their contribution to the total emission of Fe-line feature will therefore change as the temperature of a solar flare plasma changes in both space and time. This results in changes to the energies of the Fe-line feature, as defined by the energy of the peak intensity (E_p). Changes in E_p , if large enough therefore, provide a possible useful temperature diagnostic.

In Fig. 9 we show the variation of peak energy (E_p) as a function of temperature of Fe-line feature. We measured E_p for each photon spectra at a given time of a flare for which temperature was derived from continuum. A total 108 spectra from all 8 flares

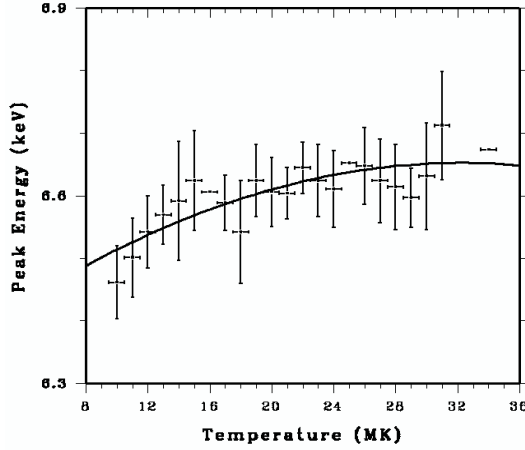


Figure 9. Variation of peak energy (E_p) of Fe-line feature as a function of temperature.

were analyzed to measure E_p . Later, in order to get better statistical confidence, we distributed the 108 E_p measured values in the interval of 1 MK according to their respective spectra temperature. For example, all E_p measurements from the spectra falling in the temperature range between 9.6 and 10.5 MK were averaged to mean value and also a standard deviation (σ) was obtained. This mean E_p value is shown with $\pm 1\sigma$ at mean temperature 10 MK. The E_p for each photon spectra of each individual flare was derived by line parameter software analysis of SOXSoft that employs Chianti code to derive plasma parameters (cf. Fig. 5). The peak energy (E_p) is the channel having peak intensity in the integrated spectra. The channel width is 0.082 keV in Si PIN spectrometer and therefore measurement of peak energy varying between 6.4 and 6.8 keV as a function of temperature could be possible with better precision. Figure 9 shows that E_p increases with temperature, which is in agreement to Phillips (2004), and Oelgoetz & Pradhan (2004).

3.1.3 Equivalent width of Fe-line features in flare plasma

The observations of the Fe-line and Fe/Ni-line features and neighbouring continuum offer a means of determining the iron abundance $A_{\text{flare}}(\text{Fe})$ and similarly the nickel abundance $A_{\text{flare}}(\text{Fe/Ni})$ during flares. The thermal plasma during flares is located in the coronal loop structures typically 10^4 km above the photosphere. On a chromospheric evaporation picture, this plasma is formed from the chromosphere and should therefore reflect the chromospheric composition. Fludra & Schmelz (1999) and Phillips *et al.* (2003) showed that elements with a variety of first ionization potential (FIP) are in ratios that are characteristic of the corona, i.e., with low-FIP (FIP ≤ 10 eV) elements enhanced by a factor of 3 or 4 but with high-FIP elements approximately the same or depleted by a factor up to 2 compared with photospheric abundance. However, element enhancement might depend upon flare intensity and duration. Thus study of a large variety of flares is important. Further Fe and Ni both are low-FIP elements and therefore SLD/SOXS observations of Fe and Fe/Ni line features in contrast to neighbourhood continuum may allow determining the abundance of Fe and Fe/Ni in flare plasma.

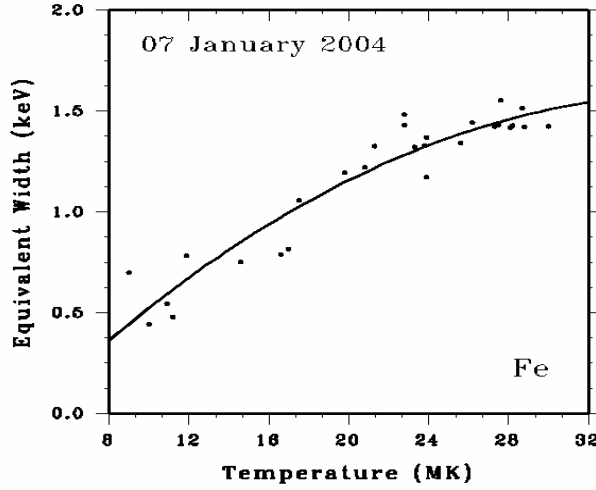


Figure 10. Variation of equivalent width (w) of Fe-line feature as a function of temperature on 07 January 2004 solar flare observed by SOXS mission.

A measure of the Fe-line feature's intensity with respect to the continuum is provided by the equivalent width (w), measured in keV, which can be determined from Si/SLD spectra. Figure 10 shows equivalent width (w) of Fe-line features as a function of temperature in 07 January 2004 flare plasma. We considered this individual flare as we have 27 integrated spectra for detailed study. It may be noted from Fig. 10 that w increases exponentially with temperature up to 20 MK and then slowly. However, the equivalent width (w), which gives an estimate of abundance of the element, is irrespective of flare location on the Sun or strength, rather it depends upon the temperature of the flare at a given time. Thus in order to get better statistical confidence we carried out a detailed study of the equivalent width (w) by analyzing all 8 M-class flares under current investigation. Like the peak energy (E_p) we derived w from each photon spectra of a given time of the flare for which temperature was measured from continuum. A total of 108 photon spectra from 8 flares were analyzed to measure the w . Later, similar to E_p , in order to get better statistical confidence, we distributed the 108 measured values of w in the interval of 1 MK according to their respective spectral temperature interval (cf. 3.1.2). For example, all w measurements from the spectra falling in the temperature range between 9.6 and 10.5 MK were averaged to a mean value and also a standard deviation (σ) was obtained. This mean w value is shown with $\pm 1\sigma$ at mean temperature 10 MK and so on. Shown in Fig. 11 is variation of the mean w with temperature, which also nevertheless shows like Fig. 10, an exponential rise of w until 20 MK and later slowly. However, it may be noted from this figure that w remains almost between 1.5 and 1.6 keV in the temperature range 28–34 MK. Temperatures more than 34 MK were not found for any flare under study so variation of w beyond 34 MK currently cannot be measured.

3.2 Ionization state in flares

Phillips (2004) showed that assumption of steady-state equilibrium in the flare plasma may not be valid due to rapid change of temperature in the rise phase of the flare.

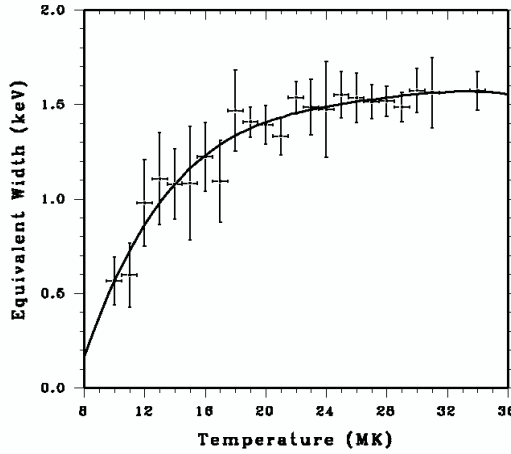


Figure 11. Equivalent width (w) as a function of temperature. Horizontal bars indicate temperature interval of ± 0.5 MK over which the mean w is derived. Vertical bars represent $\pm 1\sigma$ error in w measurements.

Therefore, if non-equilibrium conditions exist in the flares, the plasma would be expected to be in an ionizing state during the rise phase. However, the equilibrium is a good approximation unless the temperature gradient $(dT_e/dt) \geq 0.5$ MK/s (Phillips *et al.* 1974; Mewe *et al.* 1985b) if $N_e \geq 10^{10} \text{ cm}^{-3}$. Thus it is very important to derive the quantity $(dT_e/dt)/T_e$ and to observe its variation over the flare duration in order to compare with ionization and recombination time scales. We undertook this study for one flare namely 07 January 2004 out of 8 flares as we have long duration data for this flare and a total of 27 integrated photon spectra were available to derive temperature over the flare interval. In this flare we do not find (dT_e/dt) in excess of 0.08 MK/s, a factor 6 less than required for equilibrium conditions. In fact throughout the flare duration this quantity varies between -0.05 and 0.08 and therefore the flare cannot be regarded in equilibrium steady-state during rise or decay phase. Figure 12 shows three curves, viz., light curve (top panel) from Si detector in 10–20 keV energy band, T_e (middle panel), and $(dT_e/dt)/T_e$ (bottom panel) as a function of time. It may be noted from this figure that intensity of the flare and temperature variation of the flare as a function of time are almost similar. Consistent to it, the $(dT_e/dt)/T_e$ also fluctuates during rise phase from 03:55 to 04:00 UT between -0.002 and 0.0032 . During second temperature rise around 04:03 UT it coincides with intensity rise (top panel). Later part of the flare interval may be regarded as decay phase where we may note that $(dT_e/dt)/T_e$ is almost negative. Based on the measurements of temperature gradient, i.e., $dT_e/dt > 0$ from our observations during rise phase of the flare, the inverse of $(dT_e/dt)/T_e$ may be considered as ionization time, and similarly when $dT_e/dt < 0$ the inverse of $(dT_e/dt)/T_e$ may be considered as recombination time (Phillips 2004). We get ionization time of about 300 s and the whole decay phase as recombination time but in general more than 500 s.

4. Discussion

It is well established that during the flare interval the plasma is not at one temperature, rather it varies as a function of time (Feldman *et al.* 1995). However, in addition to

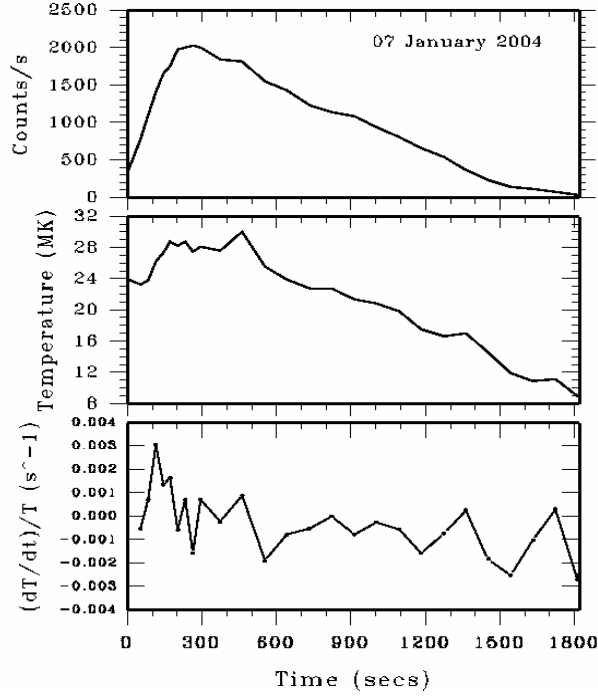


Figure 12. **Top panel:** Light curve of 07 January 2004 flare in 10–20 keV energy band as observed by Si detector of SLD/SOXS mission. **Middle panel:** Variation of temperature (T_e), and, **Bottom panel:** Variation of $(dT_e/dt)/T_e$ as a function of time on 07 January 2004 flare.

this fact, our study shows that temperature does not vary smoothly, rather fluctuates in general during the whole flare interval and rapidly in particular during rise phase (cf. Figs. 8 and 12). This fluctuation in flare plasma temperature (T_e) affects the ionization state and thereby as a consequence of it we observe variation in peak energy (E_p) and equivalent width (w) of the Fe-line emission. Our photon spectral observations from the 8 flares under study show that the minimum critical temperature required for Fe-line feature to be visible is 9 MK, which may also be seen from Figs. 9 and 10. With increase in temperature, viz., $9 < T_e < 30$ MK He-like Ca-line (3.86–3.90 keV), and He-like Fe-lines and satellite (6.4–6.7 keV) are most intense. Our Si detector begins spectral observations from 4 keV and therefore the question of observing Ca-line feature around 3.8 keV does not exist. However the observed increase in T_e unambiguously represents change in peak energy because of excitation of different principal lines of He-like Fe xxv, satellites and resonance lines in agreement to earlier calculations (Gabriel 1972; Boiko *et al.* 1978; Doschek *et al.* 1981; Feldman *et al.* 1995; Kato *et al.* 1997; Phillips 2004).

The strength of Fe-line feature above the continuum, i.e., equivalent width (w), which gives an estimate of $A_{\text{flare}}(\text{Fe})$ in terms of the coronal abundance, is also found varying over the T_e of the flare plasma. However an exponential rise in w is seen up to 20 MK and then later it slows down to remain between 15 and 1.6 keV. The equivalent width measured by us is significantly less than that calculated by Phillips (2004). This motivated us to compare with that calculated earlier by Raymond &

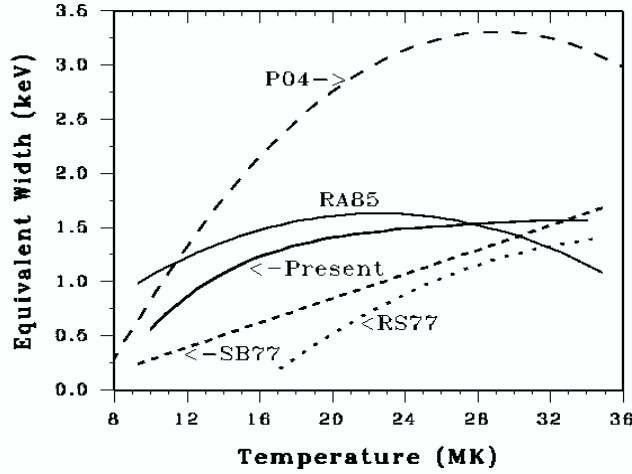


Figure 13. Comparison of our measured values of equivalent width (w) with previous results from RS77 (Raymond & Smith 1977), SB77 (Sarazin & Bahcall 1977), RA85 (Rothenflug & Arnaud 1985) and P04 (Phillips 2004).

Smith (1977); Sarazin & Bahcall (1977); Rothenflug & Arnaud (1985) and Phillips (2004) as shown in Fig. 13. It may be noted from this figure that our measured values of w are significantly higher than Raymond & Smith (1977), referred as RS77, and Sarazin & Bahcall (1977), referred as SB77, in the temperature range 10 to 30 MK. However, it comes close to SB77 around $T_e = 32$ MK but above it calculated w of SB77 moves up higher than our measured w . However, our results are slightly in better agreement with Rothenflug & Arnaud (1985), referred as RA85. Calculations of w from Phillips (2004), referred as P04, appear higher than our measurements and as well as calculations of RS77, SB77 and RA85.

Rothenflug & Arnaud (1985) calculated equivalent width (w) of the different ionic participations of Fe-line feature and showed how it varies with temperature for each individual ionic line. The w of Fe xxv line is higher than Fe xxiii, Fe xxiv and Fe xxvi up to 100 MK. Above 100 MK emissions from Fe xxvi become stronger and the total w is dominated by this emission. However, contribution to w from Fe xxii, Fe xxiii and Fe xxiv almost stops around 21, 35 and 115 MK respectively. Therefore, in the temperature range of 9–34 MK for the flares studied in this investigation, major contribution for w may be considered from these ionic emissions and Fe xxv. However, a little contribution from Fe xxvi may be considered when temperature exceeds 30 MK. The difference in calculations by earlier investigators may be due to selection of codes. Our experimental measurements of equivalent width as well as of peak energy and their variation over temperature may help to improve theoretical calculations.

5. Conclusion

The Si PIN detector of the SOXS Low Energy Detector (SLD) payload provides a unique opportunity to study the Fe-line and Fe/Ni line features in great detail. In this paper we carried out a study of Fe-line feature in order to investigate the variation

of peak energy (E_p) and equivalent width (w) as a function of temperature of the flare plasma. We found that peak energy of Fe-line feature varies from 6.4 at 9 MK temperatures to 6.7 at 30 MK. More interestingly, equivalent width (w) rise exponentially up to 20 MK and then slowly to almost reach 1.55 keV at 28 MK and beyond. We interpret the variation of both E_p and w with temperature as the changes in the ionization and recombination conditions in the flare plasma during the flare duration and as a consequence the contribution from different ionic emission lines also varies. Our measurements of w are compared with previous calculations and we found that they are close to the results of Rothenflug & Arnaud (1985). It is proposed that our measurements of w may help in improving theoretical calculations.

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EUV and Coronagraphic Observations of Coronal Mass Ejections

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Abstract. The Large Angle Spectrometric Coronagraph (LASCO) and Extreme-ultraviolet Imaging Telescope (EIT) onboard Solar and Heliospheric Observatory (SOHO) provide us with unprecedented multi-wavelength observations helping us to understand different dynamic phenomena on the Sun and in the corona. In this paper we discuss the association between post-eruptive arcades (PEAs) detected by EIT and white-light coronal mass ejections (CMEs) detected by LASCO/C2 telescope.

Key words. Sun—corona: coronal mass ejections—post-eruptive arcades—filaments—prominence.

1. Introduction

When inspecting the Sun we basically observe two different kinds of mass loss: the steady solar wind outflow and the sporadic ejection of large plasma structures which are called coronal mass ejections (CMEs). A CME is defined as an observable change in the solar coronal structure that:

- occurs over a time-scale of several minutes to hours, and
- involves the appearance and outward motion of a new discrete, bright white-light feature in the field-of-view (FOV) of a coronagraph (Hundhausen *et al.* 1984; Schwenn 1996).

CMEs propagate in the corona with speeds ranging from 50–2500 km s⁻¹ (e.g., Gopalswamy *et al.* 2003b) and are very important for space weather when they approach Earth (e.g., Webb 2004). When CMEs hit the Earth, they can produce strong geomagnetic storms interrupting satellite communication and sometimes also causing electrical power blackouts. Despite these tangible effects, the basic physical mechanism of CME initiation and their evolution is not well understood.

In order to study the initiation mechanism of CMEs, it is mandatory to precisely pin-point their source regions in the low corona and photosphere. This topic has been at the forefront of solar physics research in recent decades. Unfortunately CMEs which originate near the disk center, for which the source regions are better observed, appear

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unstructured, covering the occulter either partially or fully as they propagate parallel to the line-of-sight (LOS). Consequently they are called halos (Howard *et al.* 1982). Moreover these CMEs are often difficult to observe because of the dependence of the intensity of the Thomson scattered light on the viewing angle (Thompson *et al.* 1998; Plunkett *et al.* 1998).

After its launch in December 1995, SOHO (Solar and Heliospheric Observatory) has provided us with the observations of dynamical processes in the solar atmosphere from 1.0 to $30 R_{\odot}$ helping us study the source regions of CMEs and their evolution in the corona. Previous statistical studies showed that erupting prominences/disappearing filaments and X-ray flares (mainly long-duration: life time more than 6 hours), were found to be strongly correlated with CMEs (Sheeley 1983). Dimming regions observed in the low corona at soft X-ray and EUV wavelengths imply that the region has been evacuated and have often been found to be associated with CMEs (Biesecker *et al.* 2002 and references therein). Most of the time, dimming regions provide information about the complete base of the CME source region (Thompson *et al.* 2000).

One of the most conspicuous signatures of CMEs in the low corona is the development of arcades of bright loops on the solar disk after the CME eruption (Kahler 1977-Skylab; Hudson & Webb 1997-Yohkoh; Zarro *et al.* 1999-X-ray and EIT; Tripathi *et al.* 2004-EIT).

In this paper we present a study of the association between EUV post-eruptive arcades (PEAs) observed by the Extreme-ultraviolet Imaging Telescope (EIT; Delaboudinière *et al.* 1995) at 195 Å and white-light CMEs observed by the Large Angle and Spectrometric Coronagraph (LASCO; Brueckner *et al.* 1995) C2 aboard SOHO.

2. Observations

The EUV post-eruptive arcades (PEAs) were identified based on observations made by EIT aboard SOHO. EIT observes the Sun in four different wavelengths, namely 171 Å (Fe IX, X; 1.3×10^6 K), 195 Å (Fe XII; 1.6×10^6 K), 284 Å (Fe XV; 2.0×10^6 K) and 304 Å (He II; 8.0×10^4 K). EIT obtains one image in every 12 minutes at 195 Å which is ideal to observe the active regions and closed field regions of the quiet Sun (Moses *et al.* 1997).

The PEAs were identified using daily EIT mpeg-movies taken by EIT from 1997–2002 based on the following criteria (Tripathi 2005):

- Appearance of localized transient brightenings of large-scale loop systems ($\approx 5^\circ$) over a period of several hours.
- The large-scale loop systems which were selected had to be clearly discernible with respect to the ambient corona and observable over their full spatial extents so that their extreme end points could be located.

By definition, the criteria exclude almost all the limb events because for most of the PEAs forming at the limb the two extreme end-points cannot be uniquely defined and therefore, do not satisfy the criteria. These criteria also exclude the events which were formed after the eruption of polar-crown filaments being at higher latitude making it difficult to identify the two extreme end-points of the PEAs.

Figure 1 shows an example of a PEA identified by EIT at 195 Å on 25 January 1998 in the north-east quadrant of the solar disk. For this event the two extreme end-points

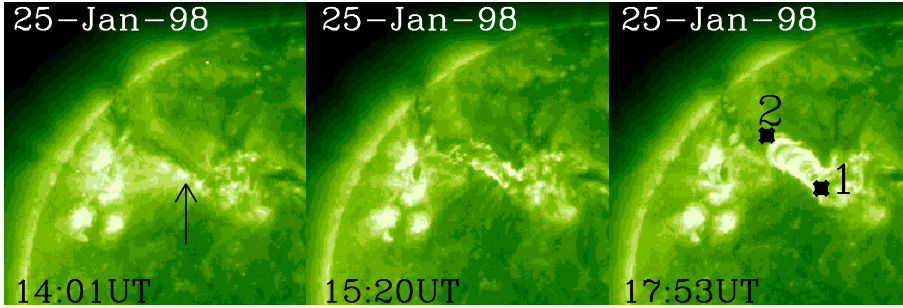


Figure 1. Images taken by EIT at 195 Å on 25 January 1998. The first two images show the eruption of a prominence (marked with an arrow in the left panel) and the consequent post-eruptive arcade. The right panel shows the post-eruptive arcade at times of its maximum evolution (emission). Point ‘1’ and ‘2’ represent the two extreme end-points of the post-eruptive arcade. In the images north points upwards and west towards the right.

of the PEA were uniquely located and are marked as ‘1’ and ‘2’ in the right panel of Fig. 1. The PEA was formed after an eruption of a prominence which is marked by an arrow in the left panel of the figure. This event has been studied in detail by Lara *et al.* (2000a, b); Gopalswamy (1999) and Gopalswamy *et al.* (1999).

Based on the above criteria, Tripathi *et al.* (2004) short-listed 236 EUV PEA events in the time period of 1997–2002. These events are tabulated and are accessible at <http://cdsweb.u-strasbg.fr/cgi-bin/qcat?J/A+A/422/337>. The table includes life-time, length, source region among other related characteristics.

Figure 2 shows the yearly variations of the total number of PEAs (blue colored bars) based on the EIT 195 Å daily mpeg-movies together with the variations of yearly averaged sunspot numbers (red colored bars). The number of identified PEAs could have been biased in early 1997 due to telemetry restriction (Subramanian & Dere 2001) and from June to December 1998 during the SOHO recovery phase after temporal disability during June–October 1998. As can be seen from Fig. 2, the total number of identified PEAs increased in the rising phase of solar activity from 1997–2000 and decreased after solar maximum in the years 2001 and 2002. Figure 2 shows a strong correlation of PEA frequency with sunspot numbers.

3. EUV post-eruptive arcades and white-light coronal mass ejections

The prime goal of this paper is to study the association of EUV PEAS observed by EIT at 195 Å with white-light CMEs detected by LASCO/C2 coronagraph. For studying the association of EUV PEAs with white-light CMEs there are basically two criteria which have to be taken into account: the spatial and temporal correlations. The spatial correlation of PEAs with CMEs was investigated by comparing the CME’s position angle (PA) as provided in the CME catalogue (Yashiro *et al.* 2002; http://cdaw.gsfc.nasa.gov/CME_list) with the location of the related PEAs on the solar disk. The PA is measured positive in the counter-clockwise direction, starting with zero degrees at the solar north. Ideally the associated PEA should lie on the same line as the radial line passing from the disk center to the center of the leading edge of the CME. However, this is always not the case and the spatial correlation alone does not always ensure that a particular CME was indeed associated with a given PEA.

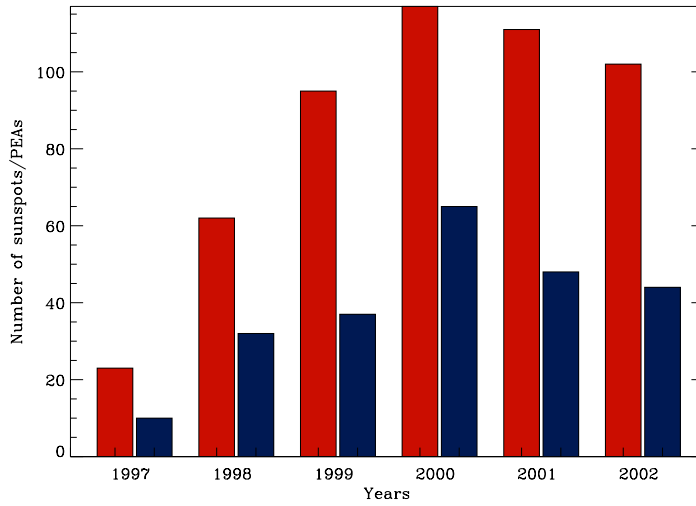


Figure 2. Averaged sunspot number (red colored bars) and total number of identified EUV post-eruptive arcades at $195 \text{ \AA}$ (blue colored bars) from 1997–2002.

Therefore, in order to support the relationships of a given PEA with CME detected by LASCO/C2, the temporal correlation of both phenomenon was further investigated by comparing the observation time of the PEA with height-time diagrams of the CME taken from the CME catalogue. The height-time diagram of the CME includes observation taken from both the LASCO/C2 (FOV $2\text{--}6 R_{\odot}$) and C3 ($4\text{--}30 R_{\odot}$) telescopes. The height-time diagrams were produced by measuring the evolution of the leading edge of the CME under the assumption that CME is evolving radially with respect to the solar limb. From the backward extrapolation of the height-time diagrams the approximated CME onset time were inferred and were compared with the approximated onset time of PEAs in EIT images.

The CME detected by LASCO/C2, associated with the PEA observed by EIT on 25 January 1998 (see Fig. 1) is displayed in Fig. 3. The CME was first detected at 15:26 UT in the LASCO/C2 field-of-view. The CME seen in LASCO/C2 was preceded by a filament eruption seen by EIT at $195 \text{ \AA}$ marked by an arrow in the left panel of Fig. 1. The estimated onset time of the CME was 14:58 UT based on the backward extrapolation of the height-time diagram. The onset time of the CME is in good agreement with the onset time of the prominence eruption observed by EIT at $195 \text{ \AA}$. The determinations of onset times from the simple extrapolations can be misleading, e.g., for those CMEs that were either accelerated or decelerated in the phase of the evolution. This can also happen in cases with Halo CMEs. The onset time of the PEA in EIT $195 \text{ \AA}$ was estimated to be about 15:02 UT. The estimated onset time for the CME and PEA is labeled in Fig. 4. This onset time of the CME is not consistent with the one reported by Gopalswamy *et al.* (1999) based on the multi-wavelength observations. This discrepancy is due to the linear backward extrapolation of the height-time diagram.

A systematic investigation was performed by Tripathi *et al.* (2004) to study the association of the entire set of PEAs identified by EIT at $195 \text{ \AA}$ with white-light CMEs detected by LASCO/C2 based on the two criteria described above.

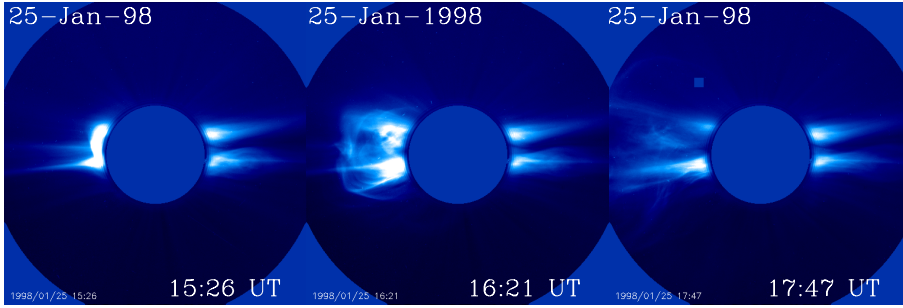


Figure 3. SOHO/LASCO/C2 white-light images showing the coronal mass ejection at the east limb associated with the post-eruptive arcade shown in Fig. 1.

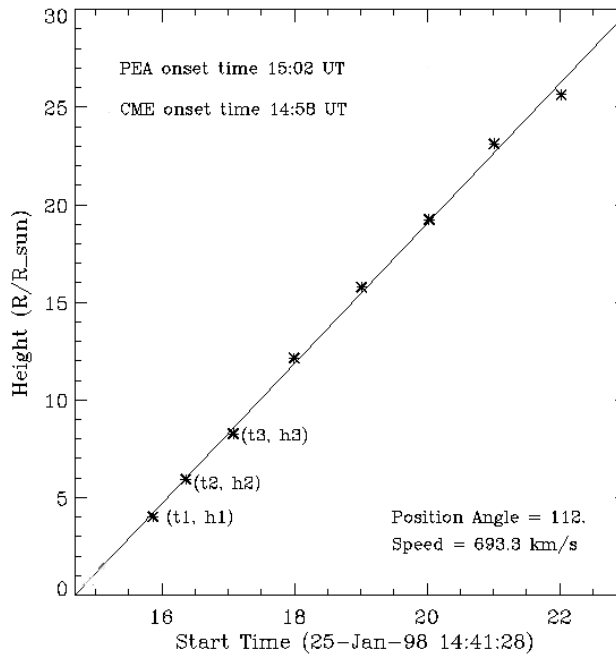


Figure 4. Height-time diagram for the leading edge of the CME observed on 25 January 1998 as provided by LASCO/C2 and C3 observations. h_1 is the height measured at time t_1 , h_2 is measured at t_2 and so on. Approximate timings of CME (triangle) and PEA (solid dot) onset is marked in the figure.

Figure 5 shows the distribution of all identified PEAs in heliographic longitude presented in bins of 10 degrees. Due to selection criteria of the events, as expected the population of identified PEAs peaks near the central meridian region where PEA events can be observed over their full spatial extent. Based on this study it was found that 92% of all the identified EUV PEAs were associated with white-light CMEs. The dashed portion of the bars represent those PEA events (other 8%) with no CMEs detected with LASCO/C2. However, from Fig. 5 it is clear that all of these events were

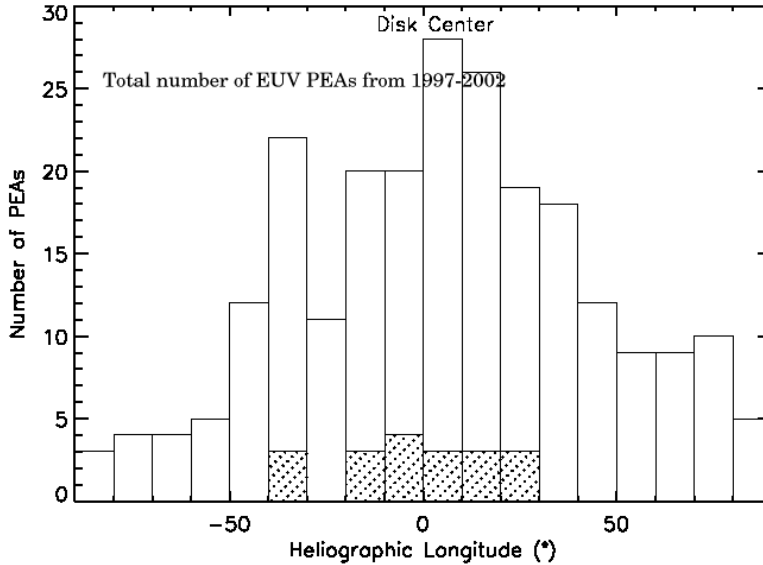


Figure 5. Distribution of PEs in heliographic longitude in bins of 10 degrees as identified in EIT 195 Å images from 1997–2002. The portions of the bars presented as the spotted areas represent those PEs for which no white-light CME had been detected by LASCO/C2 (adopted from Tripathi *et al.* (2004)).

located very near to the central meridian (in the range of $\pm 40^\circ$). The CMEs coming from the disk center are hardest to detect because of the dependence of the viewing angle on the Thomson scattered light (Thompson *et al.* 1998; Plunkett *et al.* 1998). Therefore, it seems plausible to assume that the absence of CME for some PEA events is just an observational bias and we conclude that all the PEs observed by EIT at 195 Å are associated with white-light CMEs.

Moreover, for 8% PEA events which were not associated with white-light CMEs, a comparison was made with erupting prominences and disappearing filaments based on ground-based $H\alpha$ observations. Erupting prominences and disappearing filaments are believed to be one of the clearest signatures of CMEs on the solar disk (e.g., Gopalswamy *et al.* 2003a). Taking this into account, the association between the EUV PEs and white-light CMEs increases to 98%. Therefore, we can conclude that each PEA observed at EUV wavelength (195 Å) has an associated white-light CME and so PEs can be considered as reliable disk tracers of CMEs on the solar disk even without simultaneous coronagraphic observations.

4. Summary and conclusions

Based on the observations made by EIT at 195 Å and LASCO/C2 coronagraph it was found that all the EUV PEs were associated with a white-light CME. Therefore, EUV PEs can be considered as a reliable disk tracer for the CMEs. Moreover EUV PEs can provide information about CMEs even without simultaneous coronagraph observations. These EUV PEs serve as footprints of CME source regions.

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Enigmatic Solar Wind Disappearance Events – Do We Understand Them?

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Abstract. At the Sun–Earth distance of one astronomical unit (1 AU), the solar wind is known to be strongly supersonic and super Alfvénic with Mach and Alfvén numbers being on average 12 and 9 respectively. Also, solar wind densities (average $\sim 10 \text{ cm}^{-3}$) and velocities (average $\sim 450 \text{ km s}^{-1}$) at 1 AU, are known to be inversely correlated with low velocities having higher than average densities and *vice versa*. However, on May 11 and 12 1999 the Earth was engulfed by an unusually low density ($< 0.1 \text{ cm}^{-3}$) and low velocity ($< 350 \text{ km s}^{-1}$) solar wind with an Alfvén Mach number significantly less than 1. This was a unique low-velocity, low-density, sub-Alfvénic solar wind flow which spacecraft observations have shown lasted more than 24 hours. One consequence of this extremely tenuous solar wind was a spectacular expansion of the Earth’s magnetosphere and bow shock. The expanding bow shock was observed by several spacecraft and reached record upstream distances of nearly 60 Earth radii, the lunar orbit. The event was so dramatic that it has come to be known as *the solar wind disappearance event*. Though extensive studies of this event were made by many authors in the past, it has only been recently shown that the unusual solar wind flows characterizing this event originated from a small coronal hole in the vicinity of a large active region on the Sun. These recent results have put to rest speculation that such events are associated with global phenomenon like the periodic solar polar field reversal that occurs at the maximum of each solar cycle. In this paper we revisit the 11 May 1999 event, look at other disappearance events that have occurred in the past, examine the reasons why speculations about the association of such events with global phenomena like solar polar field reversals were made and also examine the role of transient coronal holes as a possible solar source for such events.

Key words. Solar wind disappearance—polar field reversals—transient coronal holes—active regions.

1. Introduction

The origin of low-speed solar wind flows has recently been of great interest because, unlike its high speed counterpart that emanates only from large open field regions

called coronal holes, the low-speed solar wind can have different origins. Low speed solar wind is known to be associated with small mid-latitude coronal holes and it has been shown that solar wind speeds are inversely correlated with the expansion factors of magnetic flux tubes, with lower speeds coming from regions having large magnetic flux expansion factors and *vice versa* (Wang & Sheeley 1990; Sheeley *et al.* 1991). Low-speed solar wind outflows are also known to emanate from the tops of closed coronal loops in helmet streamers, from the outer boundaries of closed loop regions in active regions (Wang *et al.* 1998), and from small coronal holes in the vicinity of large active regions (Kojima *et al.* 1999). A high correlation has also been found (Nolte *et al.* 1976; Neugebauer *et al.* 1998) between solar wind speed and the size of the coronal hole from which it originates.

The extremely spectacular nature of the so-called solar wind disappearance event of 11 May 1999 has caused it to be one of the most extensively studied and reported solar wind related events in recent times. A number of observations have been reported using both space-based and ground-based instrumentation (Crooker *et al.* 2000; Farrugia *et al.* 2000; Richardson *et al.* 2000; Usmanov *et al.* 2000; Vats *et al.* 2001; Balasubramanian *et al.* 2003). Until very recently however, none of these authors were able to explain the cause or locate the solar source of the event. Many authors had speculated (Usmanov *et al.* 2000; Balasubramanian *et al.* 2003; Usmanov *et al.* 2003) that the event was somehow related to large scale solar phenomena like the periodic solar polar field reversal and had raised questions about how the solar wind could be turned off for a period of 24 hours or more. Recently, Janardhan *et al.* (2006), have been able to locate a solar source of this unusual event and have put to rest speculation about the event being associated with large scale solar phenomena like the periodic solar polar field reversal that takes place at the maximum of each solar cycle.

2. Solar wind disappearance events

As stated earlier, disappearance events are characterised by extremely low velocities and low densities. A few such events have been reported earlier, with the earliest reported case being that of the long duration density depletion in the interplanetary medium (Schwenn 1983; Gosling *et al.* 1982) which was observed by Helios 1 and Helios 2 spacecraft. The event occurred in November 1979, during solar maximum in cycle 21 and was as dramatic as the May 1999 event with proton densities decreasing steadily over an entire day from the normal value of $\sim 100 \text{ cm}^{-3}$ to less than 1 cm^{-3} , at about 0.3 AU the Helios 1 perihelion. The velocity and the magnetic field, during the same period, remained constant at 300 km s^{-1} and 35 nT, respectively. Schwenn (1983) has noted that there was no disturbance in the interplanetary medium for several days ahead of the density drop nor was there any associable solar surface feature. Usmanov *et al.* (2003) have scanned through the OMNI and ACE spacecraft databases from 1962 to 2002 and have selected those events that have densities of 0.4 cm^{-3} or less and have found a total of 18 such events. Of these 18 events, 7 including the 11 May 1999 event during Carrington Rotation 1949 (CR1949), were found to have minimum density values of 0.2 cm^{-3} or less. Table 1 shows these seven events from Usmanov *et al.* (2003). The last column in table 1 is the Alfvén Mach number and it can be seen that all the 7 events are sub-Alfvénic.

Table 1. Disappearance events with densities $\leq 0.2 \text{ cm}^{-3}$.

Date (yyyy:mm:dd)	Rotation	Day number	$\rho_{\min}$ (cm^{-3})	Alfvén Mach No. ($M_{A \min}$)
1977:10:18	1660	293	0.2	0.79
1979:07:04	1683	185	0.1	0.61
1979:07:31	1684	212	0.2	0.68
1979:11:22	1688	326	0.1	0.54
1999:05:11	1949	131	0.02	0.41
2002:03:20	1987	79	0.14	0.50
2002:05:24	1990	144	0.07	0.54

3. Spacecraft and ground based observations of the May 1999 event

The first panel (top) of Fig. 1 shows a synoptic map for CR1949 in May 1999 made using magnetograms from the MDI instrument onboard SOHO. Heliographic longitude is marked at the top while CMP dates are marked at the bottom of the map. Regions of large magnetic field strength, corresponding to active region locations are shown as black and white patches to distinguish the two magnetic polarities. The curved solid line is the magnetic neutral line. Converging black lines on the map join magnetic fields on the source surface at $2.5R_{\odot}$, derived from potential field computations (Hakamada & Kojima 1999), with their corresponding counterparts on the photosphere. The potential field lines that are marked in white correspond to fields with CMP date of 11 May 1999 and lie within the two solid, vertically oriented, parallel lines that bracket the traceback location of day-of-year 131 (DOY 131) corresponding to 11 May 1999. Also shown by thick white lines are the locations of CH boundaries inferred from HeI 10830 Å observations. The second panel shows a map of the magnetic flux expansion rates with the white, light-grey, dark-grey and black regions corresponding to flux expansion rates of > 1000 , between 100 and 1000, between 10 and 100 and between 1 and 10 respectively. The third panel shows a tomographic synoptic velocity map projected on the source surface at $2.5R_{\odot}$ for CR1949 in May 1999, obtained using Interplanetary Scintillation (IPS) data from the four-station solar wind observatory of the Solar Terrestrial Environment Laboratory (STEL), Toyokawa, Japan. The dashed white and black lines, in panel three, demarcate respectively, the boundaries of the low velocity flows ($< 400 \text{ km s}^{-1}$) and polar coronal hole boundaries. Dates of CMP are indicated at the bottom of the map with the corresponding heliographic longitudes marked at the top of the map. The path of the ACE spacecraft is indicated by a thick line along the equator. The lower panel shows ACE *in situ* measurements of proton velocities that have been mapped back along Archimedean spirals to the source surface at $2.5R_{\odot}$. The two sets of solid and dashed vertical parallel lines running across all the panels demarcate respectively, the back projected location of DOY 131 when the solar wind flow was highly non-radial and days on either side of DOY 131 when the solar wind flow was radial. The dashed vertical parallel lines therefore represent the maximum possible errors in the back projected solar source locations. Details of how the flux expansion rates are computed and a complete description of the errors in the back projection of ACE velocities from 1 AU to the source surface at $2.5R_{\odot}$ can be seen in Janardhan *et al.* (2006). It is clear from Fig. 1 (uppermost panel) that

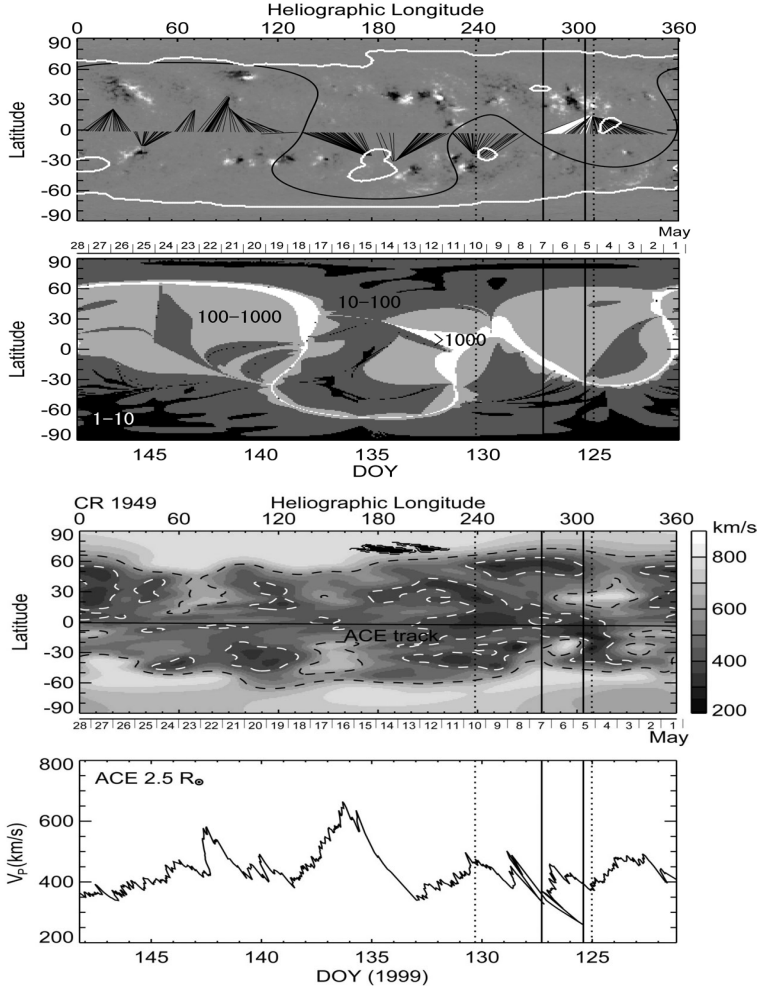


Figure 1. The uppermost panel shows a synoptic map for CR1949 in May 1999 made using magnetograms from the MDI instrument onboard SOHO. Heliographic longitude is marked at the top while CMP dates are marked at the bottom of the map. Regions of large magnetic field strength, corresponding to active region locations, are shown as black and white patches to distinguish the two magnetic polarities. The curved solid line is the magnetic neutral line. Converging black lines on the map join magnetic fields on the source surface at $2.5R_{\odot}$, derived from potential field computations with their corresponding counterparts on the photosphere. The potential field lines that are marked in white correspond to fields with CMP date of 11 May 1999. Shown by thick white lines are the locations of CH boundaries. The second panel shows a map of the magnetic flux expansion rates with the white, light-grey, dark-grey and black regions corresponding to flux expansion rates of > 1000 , between 100 and 1000, between 10 and 100 and between 1 and 10 respectively. The third panel shows a tomographic synoptic IPS velocity map projected on the source surface at $2.5R_{\odot}$ for CR1949 in May 1999. The dashed white and black lines, in panel three, demarcate respectively, the boundaries of the low velocity flows and polar coronal hole boundaries. The path of the ACE spacecraft is indicated by a thick line along the equator. The lowermost panel shows ACE in situ measurements of proton velocities that have been mapped back along Archimedean spirals to the source surface at $2.5R_{\odot}$. The two sets of solid and dashed vertical parallel lines running across all the panels demarcate respectively, the back projected location of DOY 131 when the solar wind flow was highly non-radial and days on either side of DOY 131 when the solar wind flow was radial.

the solar wind observed at 1 AU during the solar wind disappearance event originated from the vicinity of the large active region complex (AR8525) located at around 18° N, and between Heliographic longitudes 280° and 300° . Janardhan *et al.* (2006) have shown that the interplanetary magnetic field during the event was stable and unipolar, thereby implying a coronal hole origin for the solar wind flows, and have argued that the low density, low velocity flows originated from the small mid latitude coronal hole located close to the potential field lines marked in white in Fig. 1. Janardhan *et al.* (2006) have also argued that the small size of the coronal hole and the large magnetic flux expansion factors from the region could adequately explain the low velocities produced. They have further proposed an interesting method for producing the low densities, by assuming that rearrangements in CH boundaries would produce a pinch-off or separation of the solar wind outflow, thereby completely detaching the outflow from its solar source. They have argued that if such a detached outflow occurred within ~ 48 hours of its start and it continued to expand as it propagated out to 1 AU, an increase in its radius by a factor of 6–7 would lead to a decrease in densities by a factor of ~ 200 – 300 at 1 AU. Thereby, typical particle densities of approximately 20 – 30 particles cm^{-3} at 0.5 AU could be reduced to 0.1 particles cm^{-3} at 1 AU given that the typical travel time between the sun and the earth, at these low velocities, is ~ 5 days. It is important to bear in mind here that the large active region complex (AR8525) would have been located at central meridian approximately 5 days before 11 May 1999 thereby producing earth directed outflows.

4. Discussion and conclusions

4.1 Transient coronal hole

Apart from small mid-latitude coronal holes, another class of coronal holes are the so-called transient coronal holes (TCH) which were first discovered in Skylab data (Rust 1983). These TCH are short lived (≤ 2 days) regions of dimmed X-ray intensity which are sometimes observed in association with CMEs. However, coronal radiation responsible for HeI $10830 \text{ \AA}$ comes from higher altitudes than soft X-ray emission and thus represent larger spatial scales. HeI $10830 \text{ \AA}$ observations of coronal hole boundaries will therefore be blurred and cannot be used to detect the smallest CHs or TCH which are generally small in size. Even at soft X-ray wavelengths, since the corona is optically thin, foreground and background emission can prevent the detection of TCH on the limb thereby limiting TCH detections to those that are well away from the limb. Thus, detecting TCH would require special and careful processing of soft X-ray and EUV data.

In an extensive study of 19 TCH, using 9 years of data from the YOHKOH soft X-ray telescope (SXT), Kahler & Hudson (2001) have shown that TCH:

- are small in size,
- have short lifetimes;
- occur in magnetic unipolar regions trailing large active regions,
- generally occur at regions where the magnetic neutral line shows large scale curvature.

Also, the typical lifetimes of TCHs of ~ 48 hours as compared to typical travel times between the sun and 1 AU of 3–5 days for the solar wind, implies that outflows

originating at a small TCH can become totally disconnected from the source. The stable and uni-polar nature of the flows from the vicinity of AR8525 and the central meridian location of AR8525 taken together with the extensive and detailed work by Kahler & Hudson (2001) described above indicate that transient coronal holes could be an alternative solar source for explaining other such low velocity, tenuous solar wind flows seen at 1 AU. From Fig. 1 (upper panel) we can see that the solar wind flows of 11 May 1999 originated from a small area on the sun trailing AR8525 and the magnetic neutral line showed a large curvature. It is therefore not unreasonable to speculate that the other tenuous solar wind outflows listed in Table 1 could have probably originated in small TCH. The small area of the TCH associated with such an outflow would produce very low velocities (Wang & Sheeley 1990; Sheeley *et al.* 1991; Nolte *et al.* 1976; Neugebauer *et al.* 1998). The typical lifetime of a TCH of ~ 2 days as compared to solar wind travel time of ~ 5 days would imply that the low velocity flow from the TCH would have been completely detached from the solar surface approximately three days before reaching 1 AU and after propagating roughly 40% of the distance to earth orbit. As described earlier a simple expansion of this large detached low velocity flow region, as it propagated out to 1 AU, could give rise to an extremely low density cloud that engulfs the earth.

Four of the events listed in Table 1 have occurred close to the solar maximum while the remaining have occurred well within two years of the maximum when solar activity was still high. The fact that disappearance events seemed to occur at or around solar maximum when solar polar field reversals are taking place led to speculation that the large-scale restructuring of the solar magnetic fields during solar polar field reversals, is likely to be associated with density anomalies in the solar wind. The current work indicates that the connection of such events with polar field reversal periods is incidental in that the solar maximum period is dominated by large active regions and a highly deformed neutral line configuration, thereby maximizing the likelihood for the formation of small TCH (Kahler & Hudson 2001) and/or small mid latitude CH. Also, mid latitude and equatorial coronal holes are nearly absent during the solar minimum phase thereby increasing the likelihood of such events being observed only at or around solar maximum. Finally, spacecraft observations are confined to the ecliptic, another reason for observing such events at or around solar maximum because of the increased probability for the occurrence of mid-latitude and equatorial coronal holes during solar maximum.

Since coronal hole boundaries locate separatrices of coronal magnetic fields which in turn define its large-scale current systems, this work has also highlighted the need for systematic studies of the dynamics and evolution of CH boundaries. Such studies could help define coronal hole boundary structure and help in understanding boundary field connectivities. Regular and systematic observations by both ground and space based platforms will be required to identify many more such events and ground-based IPS observations will be of value in such future studies.

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Transient Perturbations and their Effects in the Heliosphere, the Geo-magnetosphere, and the Earth's Atmosphere: Space Weather Perspective

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Abstract. We discuss the effects of certain dynamic features of space environment in the heliosphere, the geo-magnetosphere, and the earth's atmosphere. In particular, transient perturbations in solar wind plasma, interplanetary magnetic field, and energetic charged particle (cosmic ray) fluxes near 1 AU in the heliosphere have been discussed. Transient variations in magnetic activity in geo-magnetosphere and solar modulation effects in the heliosphere have also been studied. Emphasis is on certain features of transient perturbations related to space weather effects. Relationships between geomagnetic storms and transient modulations in cosmic ray intensity (Forbush decreases), especially those caused by shock-associated interplanetary disturbances, have been studied in detail. We have analysed the cosmic ray, geomagnetic and interplanetary plasma/field data to understand the physical mechanisms of two phenomena namely, Forbush decrease and geomagnetic storms, and to search for precursors to Forbush decrease (and geomagnetic storms) that can be used as a signature to forecast space weather. It is shown that the use of cosmic ray records has practical application for space weather predictions. Enhanced diurnal anisotropy and intensity deficit of cosmic rays have been identified as precursors to Forbush decreases in cosmic ray intensity. It is found that precursor to smaller (less than 5%) amplitude Forbush decrease due to weaker interplanetary shock is enhanced diurnal anisotropy. However, larger amplitude (greater than 5%) Forbush decrease due to stronger interplanetary shock shows loss cone type intensity deficit as precursor in ground based intensity record. These precursors can be used as inputs for space weather forecast.

Key words. Space weather—geomagnetic storm—cosmic ray—Forbush decrease—interplanetary shock—coronal mass ejection.

1. Introduction

Space weather refers to conditions on the Sun and in the heliosphere, magnetosphere, ionosphere and thermosphere that can influence the performance and reliability of space-based and ground-based technological systems and can endanger human life or health (e.g., see Kudela *et al.* 2000). Spacecraft systems are vulnerable to space

weather through its influence on energetic charged particle and plasma population, while spacecraft electronics and aircrew are vulnerable to cosmic rays and solar particle events (Dyer and Rogers 1998). Space weather effects on the ionosphere affect electromagnetic signal propagation (Marcz 1992). On the ground, the effects of space weather perturbations are due to magnetic field effects (Lanzerotti *et al.* 1998; Pirjola 2005). Space weather predictions and forecasts, therefore, are essential for the protection of both the ground-based and space-based technological systems and other hazardous effects.

The whole space weather chain originating from the Sun, extending in the heliosphere, magnetosphere, ionosphere, lower atmosphere and to the ground is very complicated, and its full understanding, which should be aimed at for being able to forecast, and manage space weather hazards/risks, is a difficult and hard task. However, it is very important that efforts should go on in this direction.

Most of the important space weather effects are related to geomagnetic storms (GS). These geomagnetic storms are related to coronal mass ejections (CMEs) and shocks/sheath associated with them. Cosmic ray intensity, in general, is dramatically affected during the geomagnetic storms. These CMEs are potential causes of intense geomagnetic storms and decreases in cosmic ray intensity (Badrudin *et al.* 1986; Zhang & Burlaga 1988; Venkatesan & Badruddin 1990; Badruddin 1998; Cane 2000; Zhang *et al.* 2004; Kudela & Brenkus 2004; Gopalaswamy 2004).

Previous studies have observed a strong association between interplanetary CMEs and interplanetary shocks (e.g., Manoharan *et al.* 2004); and interplanetary shocks and cosmic ray intensity variations (Badrudin *et al.* 1991; Badruddin 2003); interplanetary shocks and resulting geomagnetic disturbances (Srivastava & Venkatakrishnan 2002). Therefore, to identify the shock properties that correlate better with the intensity of geomagnetic storm and to predict their arrival at Earth are the priority areas of space weather research. Attempts have been made earlier to look for the cosmic ray precursors to Forbush decreases and geomagnetic storms (Bieber and Evenson 1998; Ruffolo 1999; Munakata *et al.* 2000), but the use of cosmic ray data for space weather purposes is still in its infant stage (Kudela *et al.* 2000).

We have performed superposed epoch analysis using hourly neutron monitor data from Deep River, Climax and Calgary as a measure of cosmic ray intensity. In the method of superposed epoch analysis, the time of occurrence of certain events, e.g., CMEs, are defined as key time (zero hour/day). The data under investigation, e.g., cosmic ray intensity, is then listed sometime (hour/day) before and after the key time. In this way the data are listed in the form of a matrix in which rows represent the epochs and the columns represent time before and after the individual key time. The column averages of this matrix, which show some variations, represent the superposed epoch analysis result. As adverse space weather conditions arise due to geomagnetic storms, we have used three indices namely auroral electrojet index AE, planetary magnetic index Kp and equatorial ring current index Dst as measures of geomagnetic activity. Forbush decreases, which are transient changes in galactic cosmic ray intensity, are characterised by a fast decrease within ~ 1 day followed by a more gradual, nearly exponential recovery over a few days, have been observed continuously with neutron monitors since the 1950s. For the analysis, based on their decrease amplitude, we have divided the observed Forbush decreases in two groups (greater and less than 5%). This grouping has led us to identify the precursors to Forbush decreases and geomagnetic storms. Precursors to Forbush decreases are of practical interest as possible predictors

of space weather effects on earth, several hours or even days before the passage of a major interplanetary shock. Our results show that such efforts could be a useful input in space weather predictions.

2. Results

Figure 1 shows the superposed epoch results of hourly cosmic ray intensity data as observed at three neutron monitors before and after the onset (zero hour) of FDs of amplitudes greater than 5%. In this figure, we see an identifiable cosmic ray precursor, i.e., an intensity deficit some hours prior to the onset of decrease ('loss cone' type precursor) (see Ruffolo 1999; Manakata *et al.* 2000; Kudela *et al.* 2000). Three geomagnetic indices Dst, AE and Kp show sudden enhancement in geomagnetic activity simultaneously with the onset of Forbush decrease. Although the physical processes responsible for the two phenomena (Forbush decrease and geomagnetic storm) are thought to be different, the common cause, i.e., interplanetary shock and/or mass ejection from the Sun are responsible for both (Cane 2000; Kudela & Brenkus 2004).

In order to study the plasma/field properties of transient interplanetary events responsible for FDs of larger amplitude ($> 5\%$) and GS, we have plotted in Fig. 1 the interplanetary magnetic field strength (B), its variance (σB) and north-south component (Bz), solar wind velocity (V), their functions V.B and V.Bz.

Figure 2 shows the superposed epoch analysis plots of cosmic ray intensity variation with respect to FDs of smaller amplitude ($< 5\%$) together with the geomagnetic activity represented by indices Dst, AE and Kp. The identified precursor to FDs in this case is 'enhanced diurnal anisotropy' (see Bieber and Evenson 1998). Superposed epoch plots of interplanetary plasma and field parameters B, σB , Bz, V, VB and BzV show the simultaneously changing behaviour of these parameters. A comparison in magnitude of parameters plotted in Figs. 1 and 2 is given in Table 1.

Since the superposed epoch analysis shows only average behaviour, in Figs. 3 and 4, we have plotted hourly data of typical events, one from Group I (FD $> 5\%$) and the other from Group II (FD $< 5\%$) showing development of GS, variations in field strength, its variance and north-south component, interplanetary electric field and its dawn-to-dusk component, to show their time behaviour more clearly.

3. Discussion

A critical examination of Figs. 1 and 2 shows that there is a large difference in amplitude of cosmic ray decrease in Figs. 1 and 2. The enhanced field strength at the time of minimum intensity is the same for both sets of events, Group I and Group II (also see Table 1). The average field variance σB at the time of minimum intensity is larger for Group II events as compared to Group I events. However, the average speed of the interplanetary disturbances responsible in Group I events is larger compared to that responsible for Group II events. The product VB is almost the same in two cases. These results are consistent with the idea that stronger interplanetary shocks produce more effective modulation in cosmic ray intensity (also see Cane *et al.* 1996). Increased level of magnetic turbulence represented by increased σB , is an additional factor in causing Forbush decreases (Badrudin 2003).

Two individual events, one from each group (Figs. 3 and 4), show the variations in cosmic ray intensity, geomagnetic activity, interplanetary plasma and field parameters

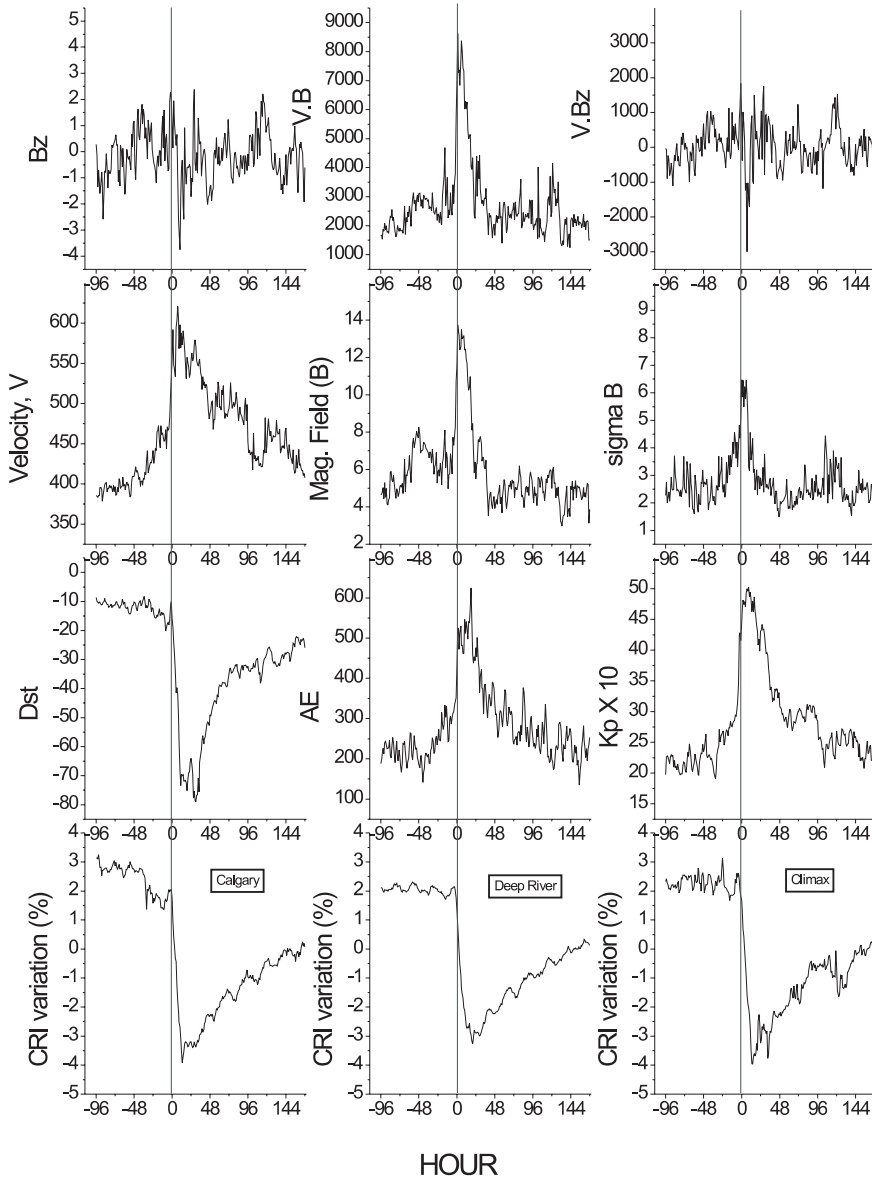


Figure 1. Superposed epoch analysis results of three neutron monitors at Calgary, Deep River and Climax, geomagnetic activity indices Dst, AE, and Kp, interplanetary plasma and field parameters, solar wind velocity (V), magnetic field (B), its variance (sigma B), north-south component (Bz), interplanetary electric field (VB), its component (VBz). Zero hour corresponds to the onset of FDs of amplitude > 5%.

before and after the arrival of disturbance. These figures also show that stronger disturbances moving with higher speed are more effective modulators of cosmic ray intensity. Further, magnetic turbulence in the interplanetary disturbance is also an important factor.

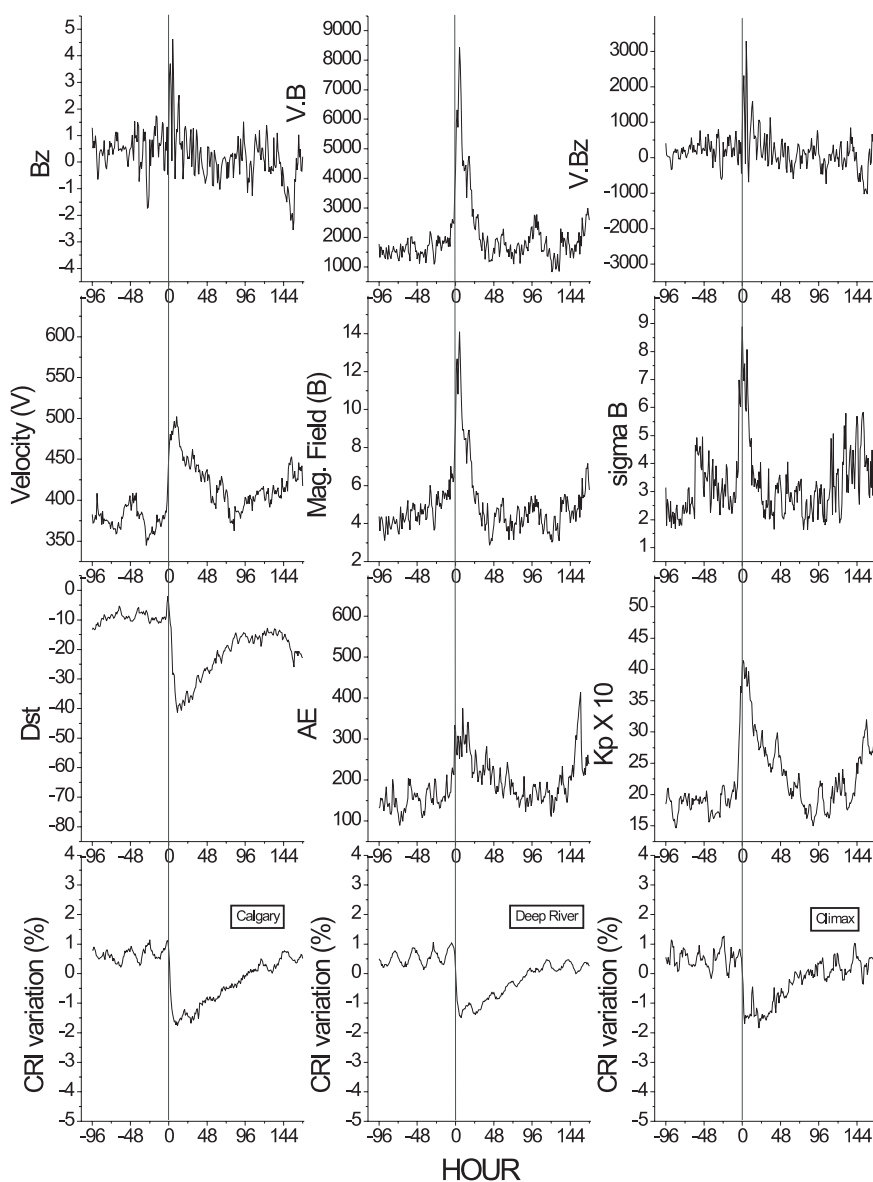


Figure 2. Superposed epoch analysis results of cosmic ray, geomagnetic and interplanetary plasma and field parameters, with respect to FDs of amplitude $< 5\%$.

The superposed epoch plots (Figs. 1 and 2) corresponding to Group I and Group II events show some noticeable features in geomagnetic activity and plasma and field parameters. On an average the magnitudes of B are the same in two cases, Bz and BzV after the passage of disturbance are larger (and negative) for Group I events than for Group II events. Moreover, the large increase in geomagnetic activity coincides with the large negative Bz periods. Similar changes are more clearly seen in individual event plots in Figs. 3 and 4. A correlation analysis between geomagnetic and solar wind

Table 1. A comparison of effects of Group I and Group II events on various parameters.

Sl. no.	Parameter	Group I	Group II
01	CRI decrease (%), Calgary NM	5.93	2.82
02	CRI decrease (%), Deep River NM	5.43	2.55
03	CRI decrease (%), Climax NM	6.58	2.68
04	Dst – Index	–78.58	–41.26
05	AE – Index	624	374
06	Kp – Index	5.04	4.16
07	Solar wind velocity, V (km/s)	623	503
08	Magnetic field, B (nT)	13.72	14.08
09	Sigma B (nT)	6.46	8.29
10	Bz (nT)	–3.74	–0.64
11	V.B	8380	8466
12	V.Bz	–2934	–680

parameters during the disturbance has been done and the best correlation is obtained with dawn-to-dusk electric field ($-VB_z$). These results support the fact that enhanced and efficient energy transfer from solar wind to the magnetosphere is due to magnetic reconnection.

After studying the interplanetary parameters responsible for geomagnetic storms and the mechanism of energy transfer responsible for GS, we have searched for the precursors that can be used for forecasting GS, some times in advance.

We have examined the precursors to Forbush decreases by analysing cosmic ray intensity recorded by three ground-based neutron detectors (see Figs. 1 and 2). The precursors to Forbush decreases are examined in association with the geomagnetic storms. The precursor to FDs of smaller amplitude ($< 5\%$) is the enhanced diurnal anisotropy, and to FDs of larger amplitude ($> 5\%$) is an intensity deficit of cosmic rays ('loss cone' type). Simultaneous analysis of solar wind, cosmic ray and geomagnetic data shows that precursors can be distinguished in terms of weaker and stronger interplanetary shocks responsible for Forbush decreases and geomagnetic storms.

A CME ejected from the Sun moving outward in the heliosphere with a speed larger than the ambient solar wind speed will produce shock/sheath region ahead of it; the shock front will hit the magnetosphere about 12–18 hours before the driving CME. As the shock reaches the ground it may produce Forbush decrease in cosmic ray intensity data and a geomagnetic storm may occur in the magnetic records almost at the same time or a few hours later. But precursors (enhanced diurnal anisotropy and cosmic ray intensity deficit) may be observed by neutron monitors many hours or even a few days before the onset of Forbush decrease and geomagnetic storm.

However, it is important that data from a number of neutron monitors, located at different latitudes and longitudes at Earth should be used to identify the precursors to FDs as all the stations may not show the precursors clearly (see Figs. 1 and 2).

It has been shown that the use of cosmic ray records in space weather research has practical application for its prediction. The observed precursors of Forbush decreases are enhanced diurnal variation and depletion of particles in a narrow loss cone. The enhanced diurnal variation of galactic cosmic ray is presumably due to an excess of

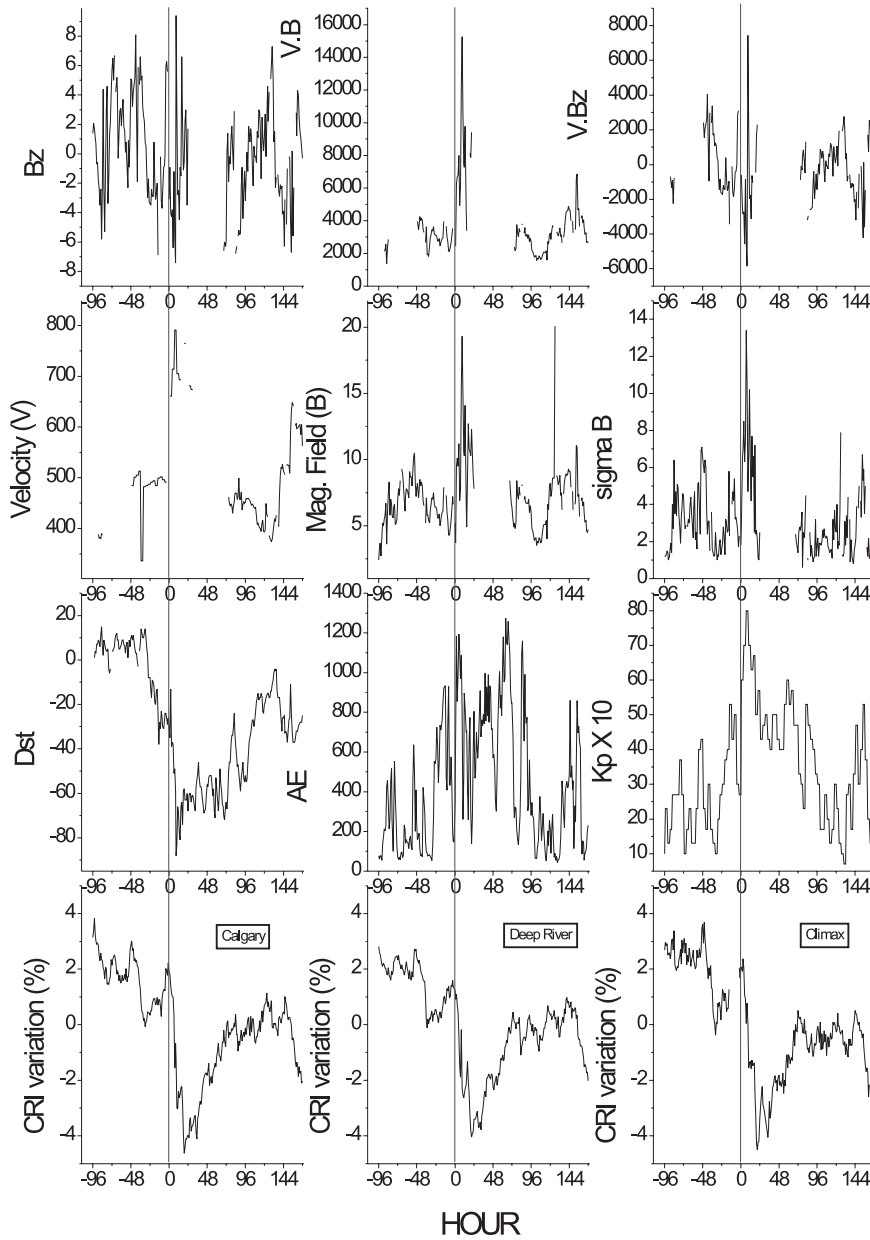


Figure 3. A typical event for Group I events ($FD > 5\%$) showing variations in various parameters before and after the onset (zero hour) of FD. Loss cone type precursory decrease is clearly seen before the onset.

particles traveling towards the Sun along the IMF when interplanetary shocks are weaker. Cosmic rays in the loss cone, i.e., along a narrow range of pitch angles presumably originate in the cosmic ray depletion region, downstream of the approaching shock when the interplanetary events are stronger.

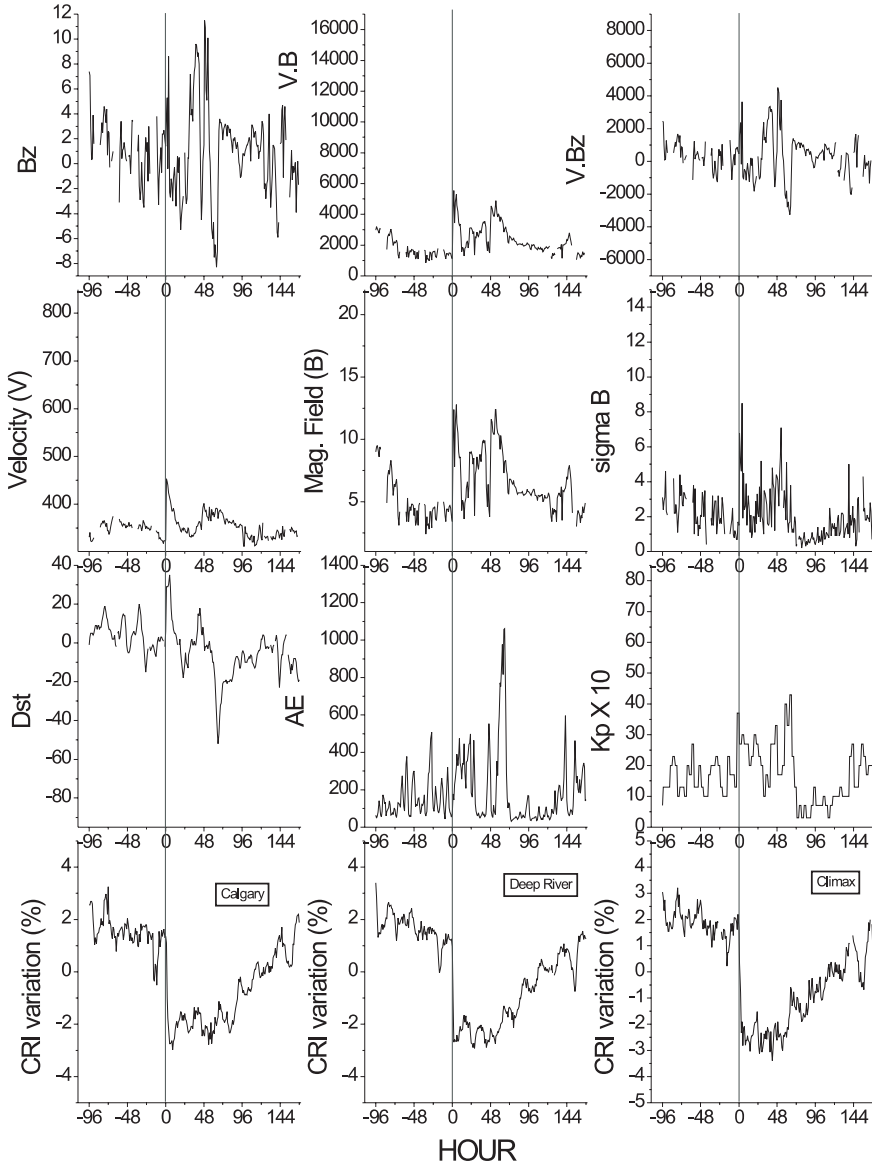


Figure 4. A typical event from Group II ($FD < 5\%$) showing cosmic ray, geomagnetic and interplanetary parameters before and after the onset (zero hour) of $FD < 5\%$, showing enhanced diurnal variation before the onset.

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Space Weather Effects of Coronal Mass Ejection

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Abstract. This paper describes the space weather effects of a major CME which was accompanied by extremely violent events on the Sun. The signatures of the event in the interplanetary medium (IPM) sensed by Ooty Radio Telescope, the solar observations by LASCO coronagraph onboard SOHO, GOES X-ray measurements, satellite measurements of the interplanetary parameters, GPS based ionospheric measurements, the geomagnetic storm parameter Dst and ground based ionosonde data are used in the study to understand the space weather effects in the different regions of the solar-terrestrial environment. The effects of this event are compared and possible explanations attempted.

Key words. Coronal mass ejection—interplanetary medium—global positioning system—total electron content—interplanetary magnetic field.

1. Introduction

The effect of coronal mass ejections and associated geomagnetic storms on space systems is called space weather. A coronal mass ejection (CME) is a large outburst of coronal magnetic field and typically 10^9 to 10^{10} tons of plasma into interplanetary space at speeds varying from 250 to 1000 km s⁻¹ (Gosling 1997). The scale and frequency of these events (several times a day to a few per week) (St. Cyr *et al.* 2000) make CMEs one of the most important contributors to space weather (McAllister *et al.* 1996; Richardson *et al.* 2000). CMEs often drive interplanetary shocks, which upon arrival on the Earth, cause geomagnetic storm. The geomagnetic storms that signal the arrival of CMEs in the near earth space pose hazards to space operations, major effects being release of trapped particles from the magnetosphere to auroral zones causing increased spacecraft charging, interference with satellite communication and surveillance systems, atmospheric heating by charged particles resulting in increased satellite drag, deterioration of magnetic torque attitude control system of satellites, etc.

During solar maximum phase, the number of such CMEs increases and many are capable of accelerating particles up to MeV energies (see special issue of *GRL*, 25(14), 1998 for a detailed review). A well observed and thoroughly investigated

CME occurred on July 14, 2000 on the Bastille day and hence called Bastille event (see special issue of *Solar Physics*, **204**, 2001 for detailed analysis of various aspects of this event). The propagation of the associated IP shock in this event down to the location of Voyager-2 at 63 AU has been identified (Burlaga *et al.* 2001). Alongwith the vigorous development of space activities, increasing interest is directed towards space weather in order to mitigate or to avoid damage by space weather calamities to technological systems caused by intense solar events. Recently Jadav *et al.* (2005) have presented the space weather aspects of a large halo CME on April 4, 2000 which appeared to be associated with 2F/C9.7 flare in AR8393. In this paper, the CME which occurred on November 4, 2001 during the violently active phase of the Sun is discussed and its space weather effects studied.

2. Observations and results

Observations of interplanetary medium using the IPS g-maps obtained from the Ooty Radio Telescope, data from SOHO/LASCO coronagraph and EIT at 171 Å; magnetic field and plasma data from the ACE/WIND satellites; Dst indices derived from ground-based magnetometer data and ionospheric observations using the GPS receivers are used to understand the complex solar-terrestrial relation during the event of November 4, 2001.

3. CME observations

LASCO and EIT onboard SOHO observed a full halo CME on November 4, 2001. The event was first observed in C2 at 1625 UT as a bright loop over the west limb; by 1650 UT the front had developed into a full halo CME (Fig. 1a and b). The CME was associated with strong X-ray flare observed by EIT at 1620 UT. This flare occurred at S17W20 quadrant in Active Region 9684 between 1603 and 1657 UT with peak emission at 1620 UT. Figure 2 shows the GOES record of X-ray flux emission which peaked at 1620 UT. A large prominence was seen erupting prior to the flare and this prominence material can be seen clearly in the LASCO images. The height-time plot from C2 and C3 observations of the position of the CME material is shown in Fig. 3. The velocity derived from height-time plot is 1868 km s^{-1} .

4. Interplanetary observations

The Ooty Radio Telescope (ORT) observes several radio sources every day and generates the g-maps which represent the excess plasma turbulence in a given direction along the line of sight to the source over the mean value. When CME material crosses the light of sight it will show up as a larger value of g in excess of unity. The ORT observations during this event are given in Fig. 4. ORT observations show significant enhancement in g-value related to the November 4, 2001 event. Radio source $1802 + 110$ showed high g-value of 2.363 on November 5, 2001 whereas for sources $1756 + 134$ g-values increased substantially to 5.118.

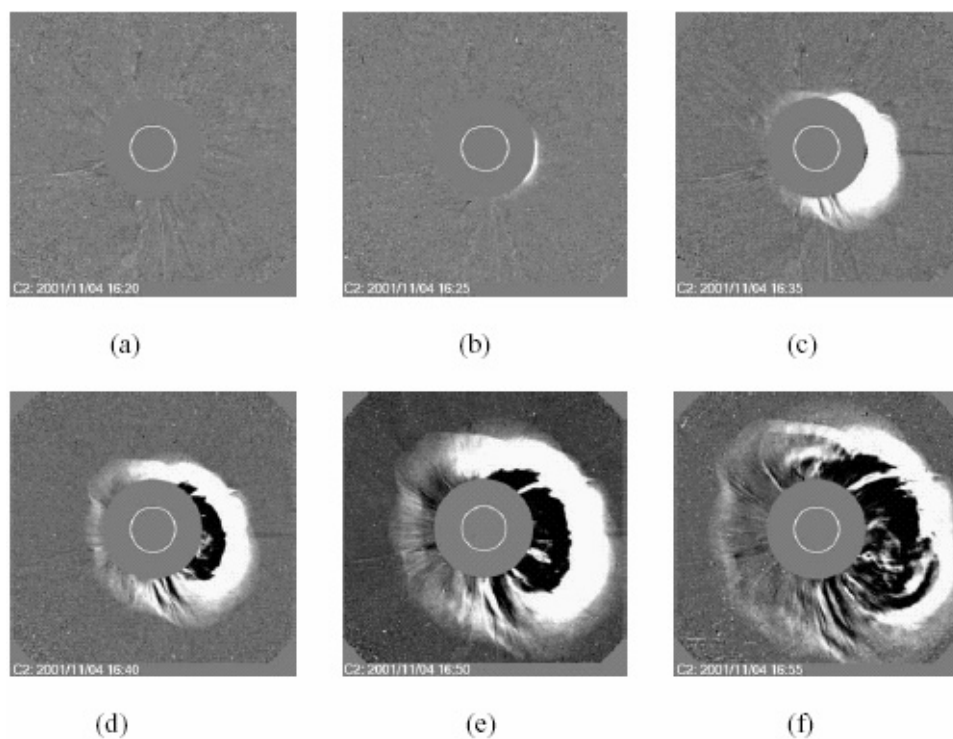


Figure 1(a–f). Image sequence of CME observed by LASCO onboard SOHO on November 4, 2001 (timings: (a) 1620 UT, (b) 1625 UT, (c) 1635 UT, (d) 1640 UT, (e) 1650 UT, and (f) 1655 UT).

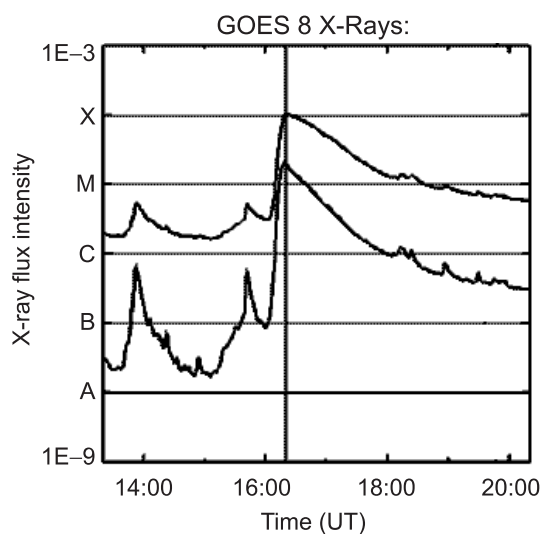


Figure 2. GOES X-ray flux data of November 4, 2001 showing an X-class flare at 1610 UT.

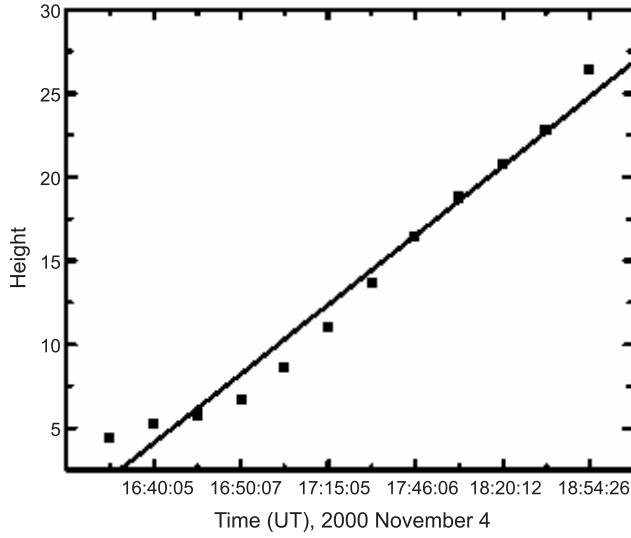


Figure 3. Height vs. time plot on November 4, 2001.

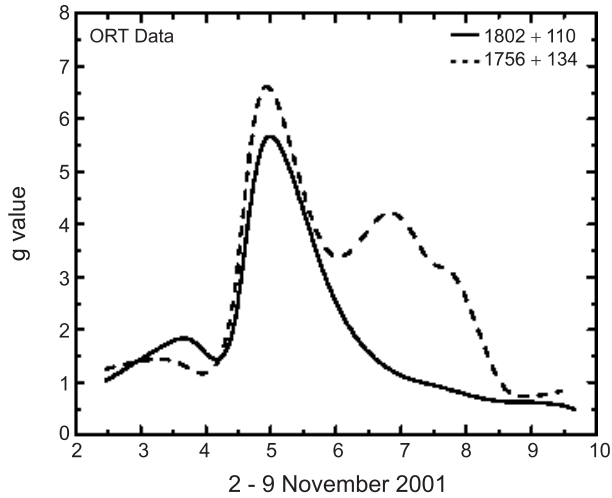


Figure 4. g-value increase observed by Ooty Radio Telescope (ORT) during November 2–9, 2001.

Measurements of interplanetary plasma and magnetic field parameters using ACE satellite during November 4–10, 2001 are shown in Fig. 5. The speed increased to $\sim 700 \text{ km s}^{-1}$ on November 6, 2001 at 0200 UT and temperature increased slightly beyond $1.35 \times 10^5 \text{ K}$. Sustained high values of density and speed were coupled with the negative B_z component of interplanetary magnetic field (IMF) which remained negative from 0600 UT on November 6, 2001 for about 5 hours. Such prolonged negative excursions of B_z are conducive for field line reconnection and the coupling of energy from solar wind to the magnetosphere which are described in the next section.

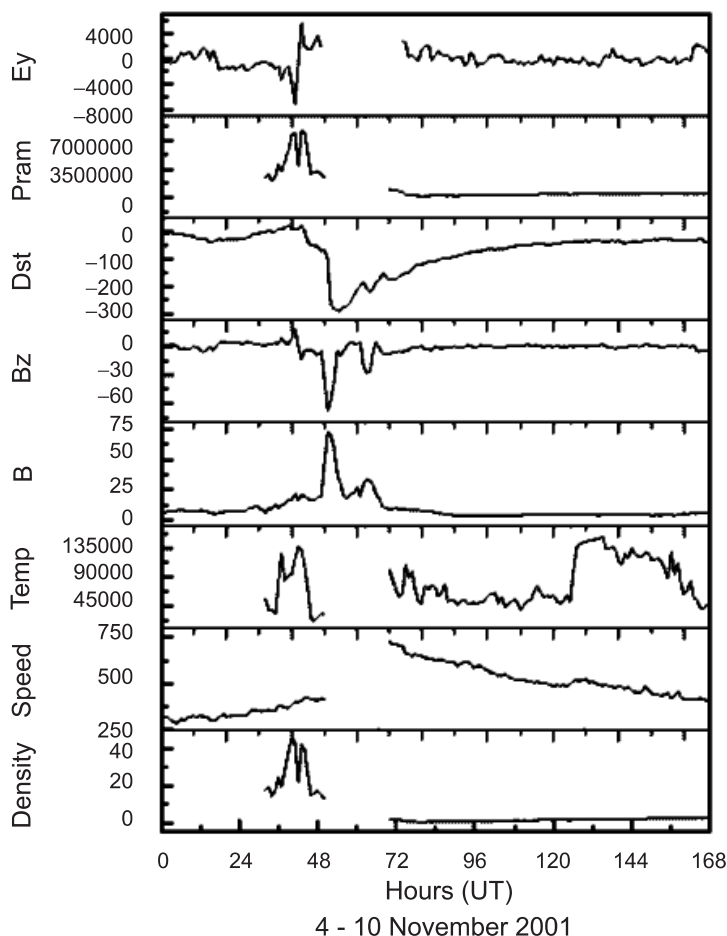


Figure 5. Interplanetary parameters at 1 AU observed by ACE satellite during November 4–10, 2001.

5. Geomagnetic and ionospheric effect

The negative excursions of B_z are well correlated with large negative excursions of the geomagnetic storm parameter Dst , which reached a minimum value of ~ -300 nT at 0600 UT on November 6, 2001 (Fig. 5). The ionospheric effects are manifested in the dual frequency TEC (TEC is the electron content in a unit column extending from the receiver to the transmitter. It is measured in TEC unit which is equal to 10^{16} electron per square meter.) measurements (Fig. 6a and b) using the ground as well as LEO satellite borne GPS receivers and the TOPEX altimeter. The composite of all these observations during 0409 UT and 0456 UT from a number of locations reveals the TEC distribution with magnetic latitude and local time (Fig. 6a). During the mid-day sector (1000–1500 local time, LT) high values of vertical TEC are observed around the magnetic equator while low values are seen around 1900 LT (see the CHAMP satellite pass). This is the typical quiet time situation. The effect of the ionospheric storm is seen in Fig. 6(b), which shows the measurements on November 6, 2001 around the same

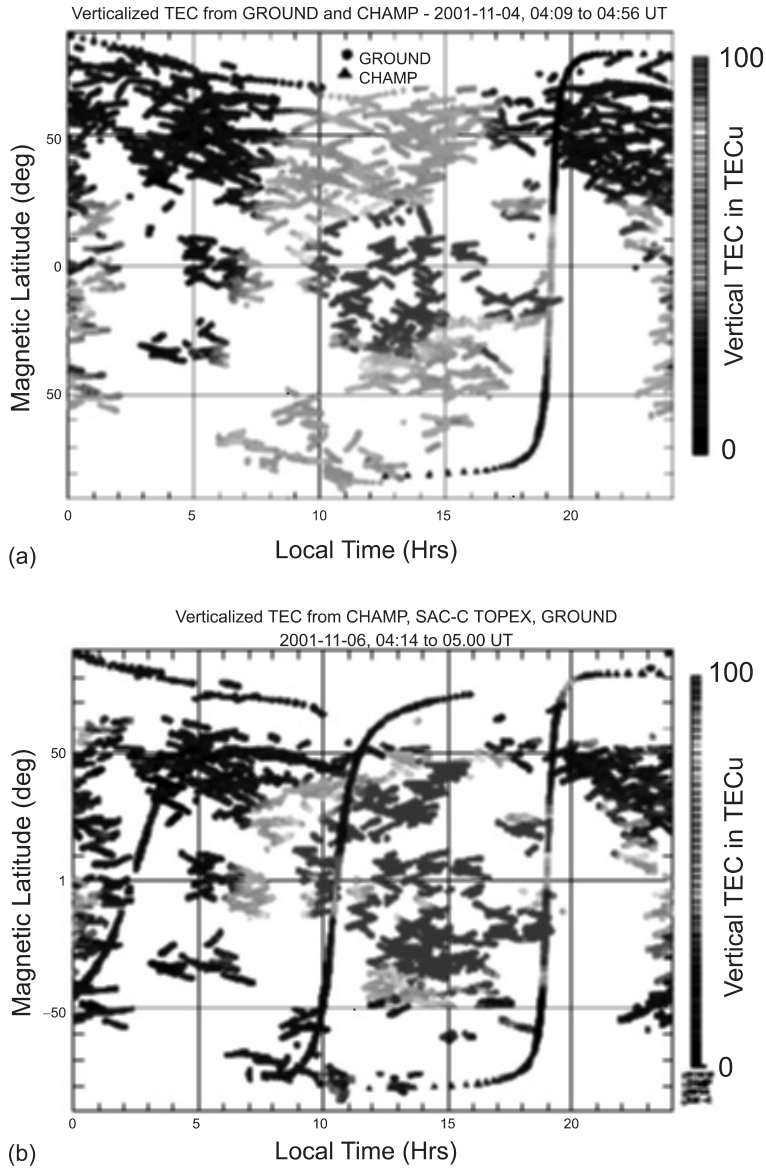


Figure 6(a–b). TEC measurements using the ground and LEO satellite-borne GPS receivers and TOPEX altimeter.

UT period but about 2 hrs after the arrival of the IP shock at 1 AU. The daytime region of high TEC has expanded to higher latitudes of $\sim 30^\circ$ N & S. In general, the TEC has increased indicating a positive ionospheric storm. The ionospheric F-layer heights at low latitudes are affected by the electric fields conveyed from the magnetosphere through the auroral region. Figure 7 shows the variations of $h'F$ (height of the bottom of the F-layer) measured at Ahmedabad. Normally at night-time, the height decreases till the F-region sunrise due to the chemical recombination, but on November 6–7, 2001

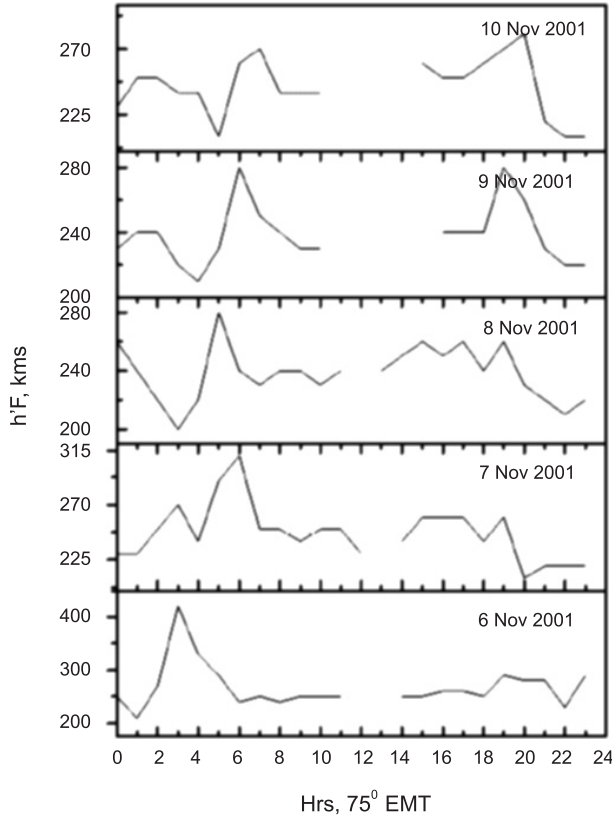


Figure 7. Variation of $h'F$ measured at Ahmedabad during November 6–10, 2001.

(during the main phase of the storm) we find an opposite trend of increasing height between 0 and 3 hrs (75° EMT). This can be attributed to the effect of the electric fields which caused the TEC redistribution shown in Fig. 6.

6. Discussion and conclusion

We have attempted to project the space weather effects associated with major CME event of November 4, 2001. The speed of this halo CME ($\sim 1868 \text{ km s}^{-1}$) with some acceleration is considered high, however the resulting geomagnetic storm on November 6, 2001 with $Dst \sim -300 \text{ nT}$ may seem relatively weak. There can be two reasons for this. One is that the period of high solar wind density ($\sim 40 \text{ particles/cc}$) was not coincident with high values ($> 750 \text{ km s}^{-1}$) of solar wind velocity. Hence, even though B_z was -60 nT , the solar wind magnetosphere coupling was rather weak. The need for positive interaction between solar wind and magnetosphere to produce geostorm was emphasized by Jadav *et al.* (2005). The other may be that, two other halo and partial CMEs were erupted from the Sun with less velocity about fifteen hours earlier. Our concerned CME may have interacted with the above two CMEs in the interplanetary medium resulting into being less powerful in producing geomagnetic storm. The detection of interplanetary disturbance (IPD) resulting from CME on November 5,

2001 by radio telescope (Fig. 4) (a day earlier than satellite) indicates the importance of interplanetary measurements in space weather studies.

The 26–34 nm Solar EUV radiation is responsible for long lived ($\sim$ hrs) TEC enhancements. Another important factor contributing to TEC enhancement at low latitudes is the penetration of magnetospheric electric fields to low latitudes. If the storm time electric fields are eastward during daytime it will enhance the $E \times B$ drift resulting in lower loss rates and increased TEC. This explains the increase in TEC at low latitudes and its possible association with solar events on November 4, 2001 around 0500 UT.

The interaction between the CME/solar wind and magnetosphere is very complex and varies from event to event. Hence, in order to make space weather prediction one has to study a large number of such events.

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Geo-effectiveness of Solar Wind Extremes

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Abstract. Examples of extreme events of solar wind and their effect on geomagnetic conditions are discussed here. It is found that there are two regimes of high speed solar wind streams with a threshold of $\sim 850 \text{ km s}^{-1}$. Geomagnetic activity enhancement rate (GAER) is defined as an average increase in Ap value per unit average increase in the peak solar wind velocity (Vp) during the stream. GAER was found to be different in the two regimes of high speed streams with +ve and –ve IMF. GAER is 0.73 and 0.53 for solar wind streams with +ve and –ve IMF respectively for the extremely high speed streams ($< 850 \text{ km s}^{-1}$). This indicates that streams above the threshold speed with +ve IMF are 1.4 times more effective in enhancing geomagnetic activity than those with –ve IMF. However, the high speed streams below the threshold with –ve IMF are 1.1 times more effective in enhancing geomagnetic activity than those with +ve IMF. The violent solar activity period (October–November 2003) of cycle 23 presents a very special case during which many severe and strong effects were seen in the environment of the Earth and other planets; however, the z-component of IMF (Bz) is mostly positive during this period. The most severe geomagnetic storm of this cycle occurred when Bz was positive.

Key words. Geomagnetic storms—geomagnetic activity—solar wind streams—interplanetary magnetic field—space weather.

1. Introduction

Solar wind is a continuous flow of hot plasma from the solar corona. This flow is supersonic and is caused due to very high coronal temperature which helps plasma overcome the gravitational attraction of the Sun. The flow is largely radial in nature and fills the entire heliosphere around the Sun. The density and speed of this flow is highly variable both in time and space. This results into the common occurrence of the plasma irregularities in the medium around planets. These irregularities give rise to a very interesting phenomenon of radio astronomy called as interplanetary scintillation. By observing the scintillation of a number of compact radio sources it is possible to estimate the solar wind speed and the spectra of plasma density in the regions of interplanetary medium (which is largely inaccessible by satellites) around the Sun. Thus the technique of interplanetary scintillation provides very valuable information about solar wind. The solar wind speed and density is also regularly monitored by

satellites. The flow of solar plasma is controlled by the energetic events, e.g., solar flares, filament eruptions, coronal mass ejections, coronal holes, etc. The volumes of data that have been collected over the last several decades have provided evidence of many extremes (both high and low type) in solar wind density and speed. Many low density and low speed events are found to correlate with the co-rotating coronal streams. The geo-environment is affected by both high and low extremes of the solar wind in a variety of ways.

Geomagnetic activity is caused by the transfer of momentum and energy from the solar wind to the magnetosphere. Rastogi (2001 and references therein) gives a historical background of the geomagnetic activity phenomena and its interesting features. Now it is an established fact that the major mechanism of energy transfer from the solar wind to the Earth's magnetosphere is magnetic reconnection (Dungey 1961). If the interplanetary magnetic field is directed opposite to the Earth's magnetic field, there is magnetic erosion on the dayside magnetosphere (by magnetic connection) and magnetic field accumulation on the nightside magneto tail region. Subsequent reconnection on the nightside leads to plasma injection at these local times. As the magneto tail plasma get injected into the nightside magnetosphere, the energetic protons drift to the west and electrons to the east, forming a current around the Earth. This current, is termed as the 'ring current', which causes a diamagnetic decrease in the Earth's magnetic field measured at near-equatorial magnetic stations. Carovillano & Siscoe (1973) have shown that the decrease in the equatorial magnetic field strength is directly related to the total energy of the ring current particles and thus it is a good measure of the energetics of the magnetic storm.

In general, the energy and momentum transfer occurs in response to the reversal of IMF or when B_z turns southward and remains so. When IMF possesses a southward component, reconnection between oppositely directed IMF and the magnetospheric fields occurs on the dayside magnetopause (Aubry & McPherron 1971). Snyder *et al.* (1963) showed a possible link between interplanetary solar wind speed V and geomagnetic index K_p , such a relationship has been examined by many workers and found to be rather loose.

Geomagnetic storms can occur due to several interplanetary features, including magnetic clouds, but it is only when the magnetic field of the interplanetary feature engulfing the Earth has a strong southward component (Wang *et al.* 2003). The interplanetary causes of geomagnetic storms have been extensively studied (Tsurutani *et al.* (1992) and references therein). The Earth-directed solar wind speed and southward component of interplanetary magnetic fields are of utmost importance in enhancing geomagnetic disturbances. In this article two distinct cases of solar wind extremes are reported for their geo-effectiveness in the following sections;

- high speed streams which were found to be associated with solar flares and coronal holes are discussed for their geo-effectiveness, and
- interesting period of most violent solar activity during October–November 2003.

The period October–November 2003 was very important for space weather investigations and to understand various geophysical phenomena. A special issue of *Geophysical Research Letters* 32(3), 2005, published a series of research papers covering various aspects of solar terrestrial relationship during this period.

2. High speed streams

A high speed stream is identified by a large increase ($\geq 100 \text{ km s}^{-1}$) in solar wind in a day and lasting for several days. Using these criteria Lindblad & Lunstedt 1981, 1983 and Lindblad *et al.* 1989 made catalogues of high speed solar wind streams from the observations made by near Earth probes and spacecraft for the period 1964 to 1982. These streams were associated with coronal holes and solar flares. There were 637 high speed streams from which 117 were associated or generated by solar flares. Vats (1992a) found that the total duration of high speed streams and its product with mean enhancement in speed correlates very well with solar activity. Mavromichalaki *et al.* (1988) prepared a catalogue of 423 well-defined high speed streams from the solar wind observations during 1972–1984. Vats (1992b) used this catalogue and found that

- flare associated streams created more severe storm (~ 2.7 times more Ap enhancement) than coronal hole associated streams and
- coronal hole associated streams caused a geomagnetic storm of longer duration than flare associated streams.

The temporal plots of Ap (continuous curve) and Dst (dotted curve) are shown in Figs. 1 and 2 for July 1975 and July 1980 respectively. There are straight lines marked on them as CS +, CS –, FGS + and FGS –. CS and FGS stand for coronal hole associated streams and flare associated or generated streams respectively. The signs + and – indicate the polarity of Interplanetary Magnetic Field (IMF). In these examples, it is interesting to note that both CS and FGS with +ve and –ve IMF produce enhancement in geomagnetic activity.

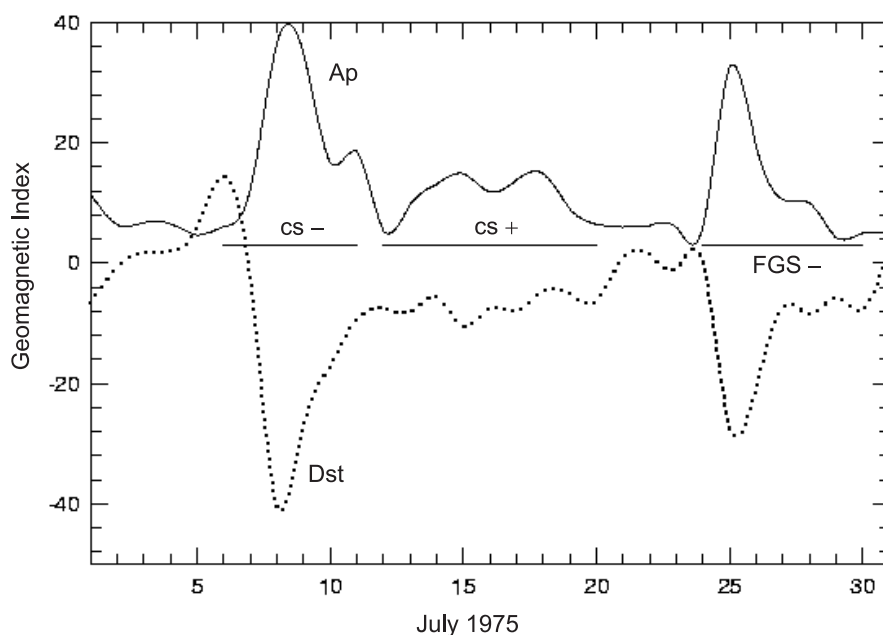


Figure 1. Temporal variation of geomagnetic indices Ap and Dst during July 1975, the straight lines marked with CS–, CS+ and FGS– show the duration of high speed streams.

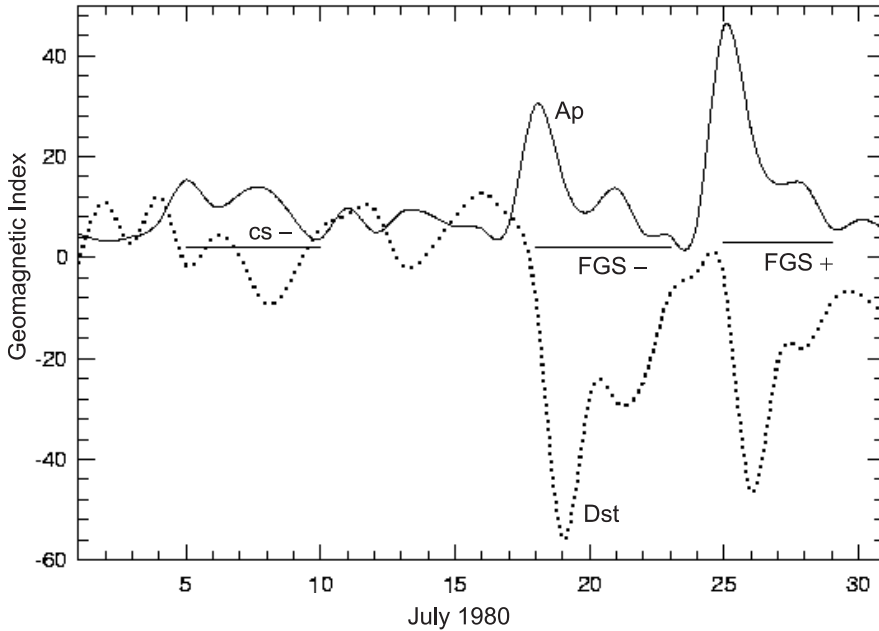


Figure 2. Same as Fig. 1 for July 1980, the straight lines marked with CS–, FGS– and FGS+ show the duration of high speed streams.

All the 423 high speed streams were grouped in two sets with +ve and –ve polarity from Mavromichalaki *et al.* (1988) catalogue. These sets were further divided in subgroups with speed increments of 100 km s^{-1} . For each of these subgroups average enhancement in Ap values were calculated. To examine the effect of increase of peak velocity V_p of the streams on the enhancement of Ap, average V_p was calculated for the subgroups. The average Ap enhancement for each subgroup was plotted with average peak velocities V_p of the same subgroup. These curves for streams with –ve IMF (dotted curve) and for streams with +ve IMF (continuous curve) are shown in Fig. 3. From this (Fig. 3), it is clear that the streams with +ve and –ve IMF are both effective for the enhancement of magnetic activity. These curves for +ve and –ve IMF have two slopes, one for $V_p < 850 \text{ km s}^{-1}$ and other one for $V_p > 850 \text{ km s}^{-1}$. Thus, the high speed streams with velocity $< 850 \text{ km s}^{-1}$ can be termed as extremely high speed streams. It is possible to define ‘Geomagnetic activity enhancement rate’ (GAER) as below:

$$\text{GAER} = \frac{\text{Increase of average Ap value}}{\text{Increase of average } V_p}.$$

GAER for streams with +ve IMF are 0.08 and 0.73 for $V_p < 850 \text{ km s}^{-1}$ and $V_p > 850 \text{ km s}^{-1}$ respectively. Similarly GAER for streams with –IMF are 0.09 and 0.53 in the two regimes of V_p . This indicates that the extremely high speed streams ($< 850 \text{ km s}^{-1}$.) with IMF +ve and –ve are more effective in enhancing geomagnetic activity than the high speed streams with lower V_p . The ratio of GAER for the extremely high speed streams with +ve and –ve IMF is 1.4 which indicates that streams with +ve IMF in this regime ($V_p < 850 \text{ km s}^{-1}$) are relatively more effective

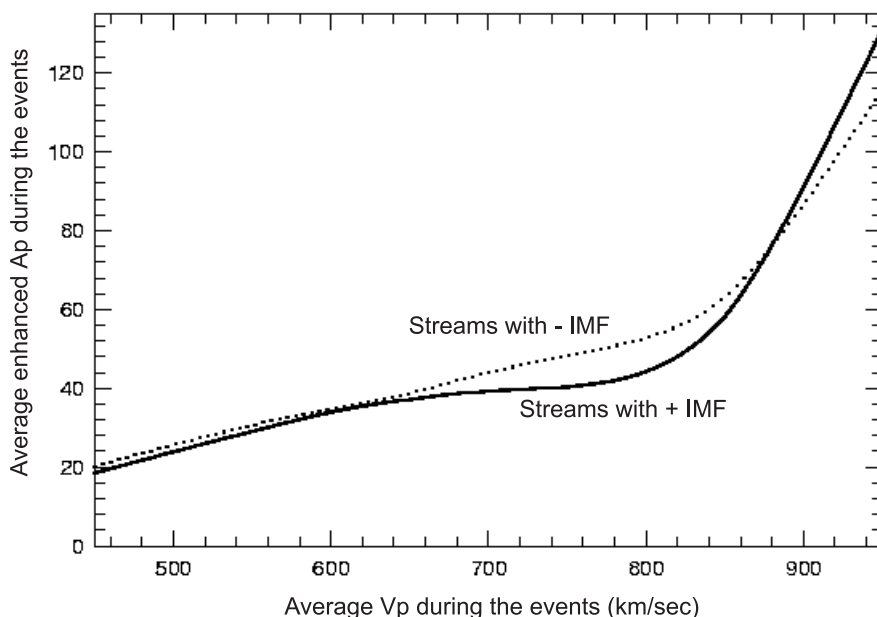


Figure 3. Ap enhancement as a function of Vp for the streams with + and –IMF.

to enhance geomagnetic activity than those with –ve IMF. However, for high speed streams $V_p < 850 \text{ km s}^{-1}$ the value of GAER 0.08 and 0.09 for streams with +ve and –ve IMF respectively. This indicates that the streams with – IMF are slightly (~ 1.1 times) more effective in enhancing the geomagnetic activity.

3. Violent solar period

During October–November 2003, three solar active regions produced a series of eruptions. These eruptions were extreme in terms of their origin and terrestrial consequence (Gopalswamy *et al.* 2005). Kane (2005) remarked that the largest Dst value was -589 nT on 13 March 1989 with velocity (V) only $\sim 550\text{--}800 \text{ km s}^{-1}$. The next largest Dst values were near -400 nT , but their V values were vastly different (Halloween events, 29 October 2003, Dst -401 nT , $V \sim 2000 \text{ km s}^{-1}$ and 20 November 2003, Dst -472 nT , $V \sim 700 \text{ km s}^{-1}$). Figure 4 shows the solar X-ray observations of GOES satellite during the period of October–November 2003. This period had a large number of solar flares, three flares on October 28, 29 and November 4 were of class X17.2, X10 and X28 respectively. The X28 was the largest solar X-ray flare observed and recorded so far. All the three of these flares are one among the 14 most powerful flares observed in the last three decades. Thus the Sun was most active and violent during October–November 2003. Bhardwaj *et al.* (2005), reported that during November 26–29, 2003 XMM-Newton observed soft (0.2–2 keV) X-ray emission from Jupiter for 69 hours. The simultaneous light curves of Jovian and solar X-rays show similar day-to-day variability. A large solar X-ray flare occurring on the Jupiter-facing side of the Sun was found to have a corresponding feature in the Jovian X-rays. Thus X-ray emission from Jovian low-latitudes appears to be solar X-rays scattered

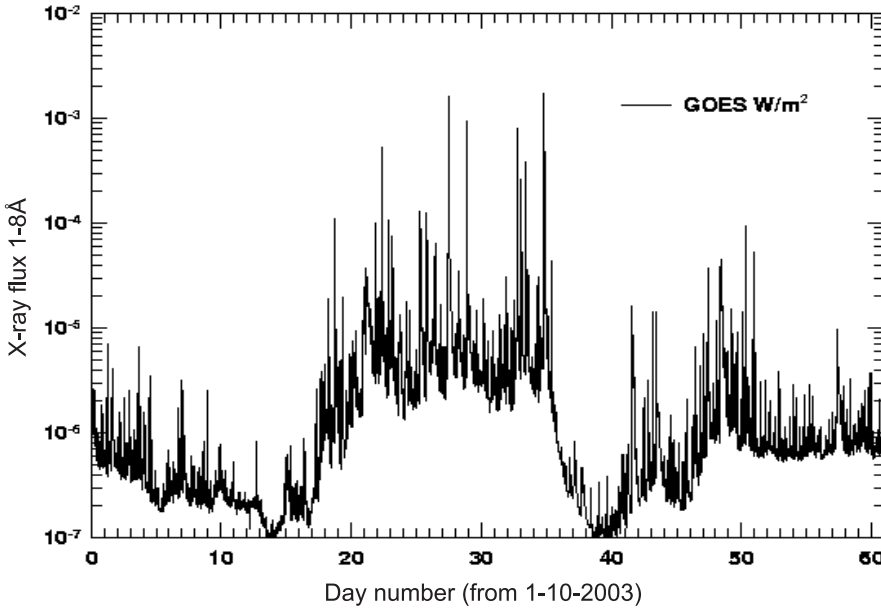


Figure 4. Temporal variation of solar X-ray flux during October–November 2003.

from the planet's upper atmosphere. Here we report only the geomagnetic effects due to violent solar activity period by way of investigating the temporal variation of Ap and Dst and the associated IMF conditions.

The temporal variation of Ap and Dst is shown in Fig. 5. The Dst values in Fig. 5 are negative during the entire period, indicating that this period was geo-magnetically disturbed. The enhancements in Ap and the depletions in Dst were simultaneous and at the same time very large too. Almost everyday there was at least one geomagnetic storm; ranging from weak to very strong one. October 28–31 and November 20, 2003 were the most violent in terms of geomagnetic activity. Gopalswamy *et al.* (2005) reported that the largest geomagnetic storm of the solar cycle 23 occurred on November 20, 2003 with a Dst index of -472 nT, due to a coronal mass ejection (CME) from active region 0501. The CME near the Sun had a sky-plane speed of ~ 1660 km s $^{-1}$, but the associated magnetic cloud (MC) arrived with a speed of only 730 km s $^{-1}$.

The temporal variation of north–south component of IMF (Bz) during October–November 2003 is shown in Fig. 6. This figure shows that Bz during October–November 2003 has more +ve excursions than –ve. The period during which Bz was +ve is larger than that for –ve. The magnitude of all the negative values of Bz during this period were less than 10 whereas positive values of Bz went up to as high as < 50 nT. Comparing the temporal variations shown in Figs. 5 and 6, it is clear that:

- whenever there is a severe geomagnetic activity indicated either by the enhancement in Ap value or by a corresponding decrease in Dst value, the Bz is positive, and
- on November 20, 2003, Ap and Bz are most positive and the Dst is most negative.

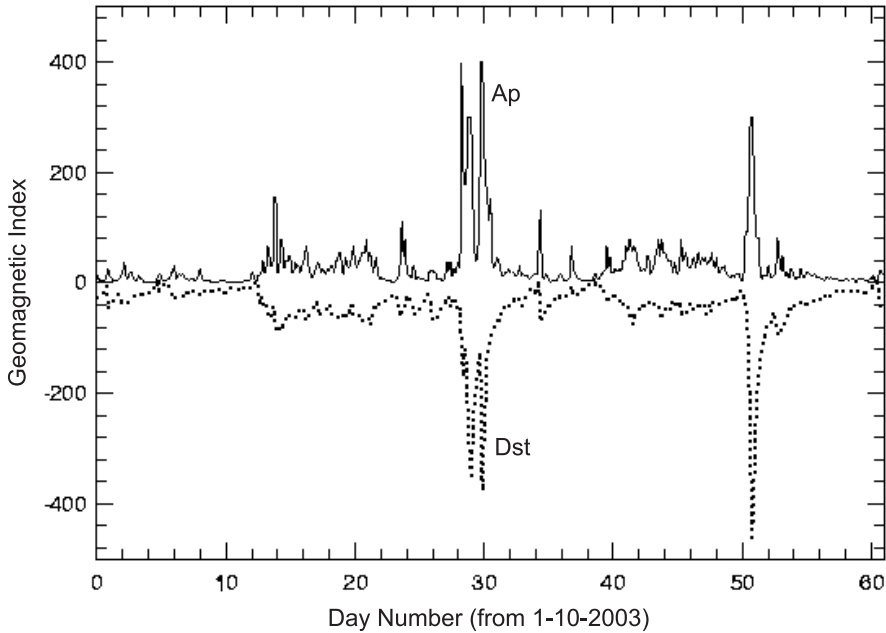


Figure 5. Temporal variation of Ap and Dst during October–November 2003.

Thus, the most violent geomagnetic storm took place when Bz is highly +ve. This is contrary to the common belief that the geomagnetic activity is always associated with negative Bz. Another associated space weather complexity was observed in the form of first day time auroral records during this violent solar period (Pallamraju and Chakrabarti 2005). They reported high spectral resolution measurements of OI 6300 Å emissions during daytime on 30 October 2003 from Boston. These observations revealed prolonged auroral activity from 1415–1900 h LT. Intense enhancement in brightness of 38 K Rayleighs was observed at 1509 h LT, which was six times larger than the normal daytime emission rates for that day. During this event the solar zenith angle was 74 Å and the solar background continuum was 4×10^6 Rayleighs Å⁻¹. During this period of auroral observations Bz was relatively high and +ve.

4. Conclusion and discussion

The examples of high speed streams associated with solar flares (FGS) and coronal holes (CS) with +ve and –ve IMF and the corresponding Ap and Dst variation indicate clearly that the high speed streams FGS and CS both are effective in enhancing geomagnetic activity. There were geomagnetic storms with both +ve and –ve IMF. The geomagnetic activity enhancement rate (GAER) is different in the two regimes of high speed streams. From the investigations of coupling between interplanetary parameters and geomagnetic activity, Ballatore (2002) found that the correlation coefficients obtained for data points corresponding to solar wind slower than 550 km s⁻¹ are equal or slightly higher than the global correlations. The observations showed that generally lower correlation coefficients for faster solar wind speeds < 550 km s⁻¹.

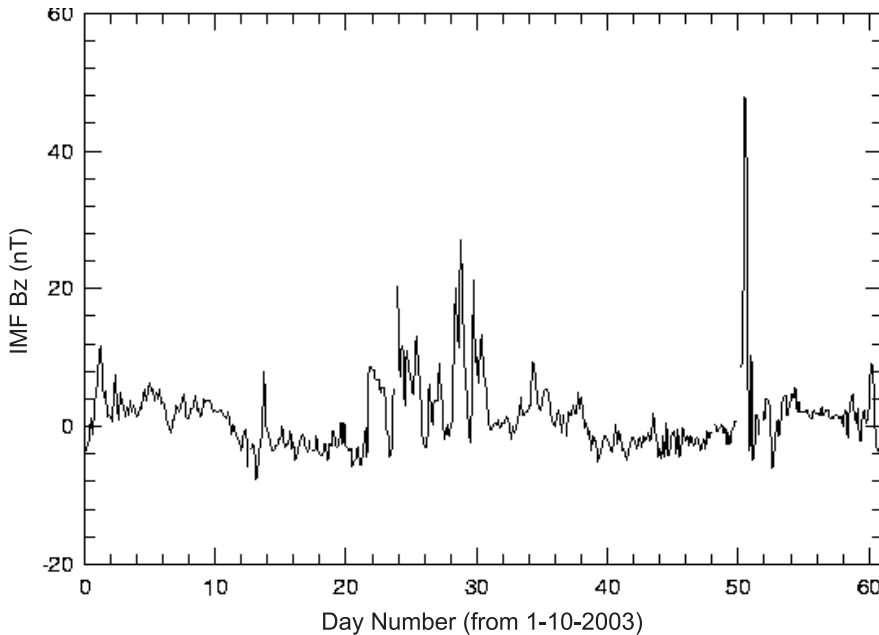


Figure 6. Temporal variation of Bz during October–November 2003.

These results are quite in contrast with the present studies. However, the one similarity between these two studies is that the relationship between the interplanetary medium and the geomagnetic activity are found to be different during periods of solar wind faster or slower than a threshold speed. The threshold speed in the present analysis was found to be 850 km s^{-1} whereas; in Ballatore (2002), it was considerably less (550 km s^{-1}). The processes responsible for the coupling between the interplanetary medium and the magnetosphere could certainly be different during different regimes of solar wind speed. Fenrich & Luhmann (1998) studied 29 magnetic cloud events and found that 40–45% of magnetic clouds, independent of polarity, are followed by fast streams which compress the tail end of the cloud. The compression results in an increase in the solar wind density and in 64% of the cases an increase in magnetic field strength also. Such tail end compression may have a significant effect upon geomagnetic activity if the magnetic cloud is of N–S polarity. The observations of violent solar period are unique in that major geomagnetic disturbances and daytime auroral emissions are associated with +ve Bz. In the literature Bz –ve is usually termed as a necessary condition for the creation of severe geomagnetic disturbance and intense aurora. The examples presented here are not in agreement with this belief. Thus more careful analysis is needed to ascertain the role of polarity of Bz in transferring the solar wind energy to magnetosphere and beyond.

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<http://spidr.ngdc.noaa.gov/spidr/>. The research at Physical Research Laboratory is supported by Department of Space, Government of India.

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The Challenge of Predicting the Occurrence of Intense Storms

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Abstract. Geomagnetic super-storms of October and November 2003 are compared in order to identify solar and interplanetary variables that influence the magnitude of geomagnetic storms. Although these super-storms ($D_{ST} < -300$ nT) are associated with high speed CMEs, their D_{ST} indices show large variation. The most intense storm of November 20, 2003 ($D_{ST} \sim -472$ nT) had its source in a comparatively small active region and was associated with a relatively weaker, M-class flare, while the others had their origins in large active regions and were associated with strong X-class flares. An attempt has been made to implement a logistic regression model for the prediction of the occurrence of intense/super-intense geomagnetic storms. The model parameters (regression coefficients) were estimated from a training data-set extracted from a data-set of 64 geo-effective CMEs observed during 1996–2002. The results indicate that logistic regression models can be effectively used for predicting the occurrence of major geomagnetic storms from a set of solar and interplanetary factors. The model validation shows that 100% of the intense storms (-200 nT $< D_{ST} < -100$ nT) and only 50% of the super-intense ($D_{ST} < -200$ nT) storms could be correctly predicted.

Key words. Sun: coronal mass ejection—geomagnetic storms—space weather.

1. Introduction

Space weather prediction involves forecasting of the magnitude and the time of the commencement of a geomagnetic storm, based on solar and interplanetary observations. Most of the currently used prediction schemes are based on the formula of Burton *et al.* (1975). They are generally reliable, even though they depend solely on interplanetary (IP) parameters, viz., the solar wind speed and the southward component of the interplanetary magnetic field. However, when solar inputs are used for prediction, one encounters the problem of ‘false alarms’ (the predicted events never occur) and of ‘missing alarms’ (no obvious solar signatures of the geomagnetic storm) as also reported by Schwenn *et al.* (2005). Although the presently available prediction schemes accurately predict the D_{ST} index of the geomagnetic storm, their prior warnings are only a few hours ahead of the commencement of the geomagnetic storm. This is because they are based on *in-situ* properties of the solar wind that can only be measured close to the Earth. For example, the ACE spacecraft measurements give

about 30 to 60 minutes of warning time as it measures the solar wind properties at the L1 point. A longer warning time requires a solar wind monitor further upstream (McPherron *et al.* 2004). For advance prediction of the arrival time of a CME at the Earth, it is necessary to

- identify the solar origins of geomagnetic storms and
- understand their relation with interplanetary parameters, viz., solar wind speed and the B_z component of the IMF that are responsible for producing super-storms.

To achieve this goal, it is also important to develop a prediction scheme based on both solar and IP properties of geo-effective CMEs.

In the present paper, key solar and IP parameters which may have been responsible for producing super-storms ($D_{ST} < -300$ nT) during October–November 2003 are described. Using these parameters for a larger data-set of 64 geo-effective CMEs recorded during 1996–2002, a logistic regression model was developed to forecast the occurrence of an intense or super-intense storm. For the development of the prediction model, the geomagnetic storms with $D_{ST} < -200$ nT were considered as ‘super-intense’ while the storms with -200 nT $< D_{ST} < -100$ nT were considered as ‘intense’.

2. Super-intense storms of the current solar cycle

During October–November 2003, several powerful CMEs occurred on the Sun which led to intense/super-intense geomagnetic storms (Gopalswamy *et al.* 2005a, b; Srivastava 2005a). Out of these, three super-intense storms that were associated with CMEs and active region flares are described here in Table 1. Although these CMEs emanated at very high speeds for instance, ~ 1175 km s $^{-1}$ (November 18, 2003) and 2125 km s $^{-1}$ (October 28, 2003), the source active regions of these CMEs showed a remarkable difference in area. The November 18 CME originated from a smaller region and was associated with a smaller and weaker flare (M3.9, 2N) as compared to the CMEs of October 28 and 29, both of which originated from a larger active region (Figs. 1 and 2). All the three geo-effective CMEs gave rise to strong interplanetary

Table 1. Source regions of super-storms during October–November 2003.

Date	October 28, 2003	October 29, 2003	November 18, 2003
Source flare	X17.2 4B flare in NOAA AR 10486 at 10:50 UT	X10.2 2B flare in NOAA AR 10486 at 20:50 UT	M3.9, 2N flare in NOAA AR 10501 at 8:30 UT
Active region area (corrected hemispherical area)	$2500 \times 10^{-6} A_{\odot}$	$2500 \times 10^{-6} A_{\odot}$	$370 \times 10^{-6} A_{\odot}$
Speed in LASCO-C2	2125 km s $^{-1}$	1950 km s $^{-1}$	1175 km s $^{-1}$
IP shock speed	~ 2000 km s $^{-1}$	1000 km s $^{-1}$	~ 730 km s $^{-1}$
D_{ST} index	-363 nT	-400 nT	-472 nT

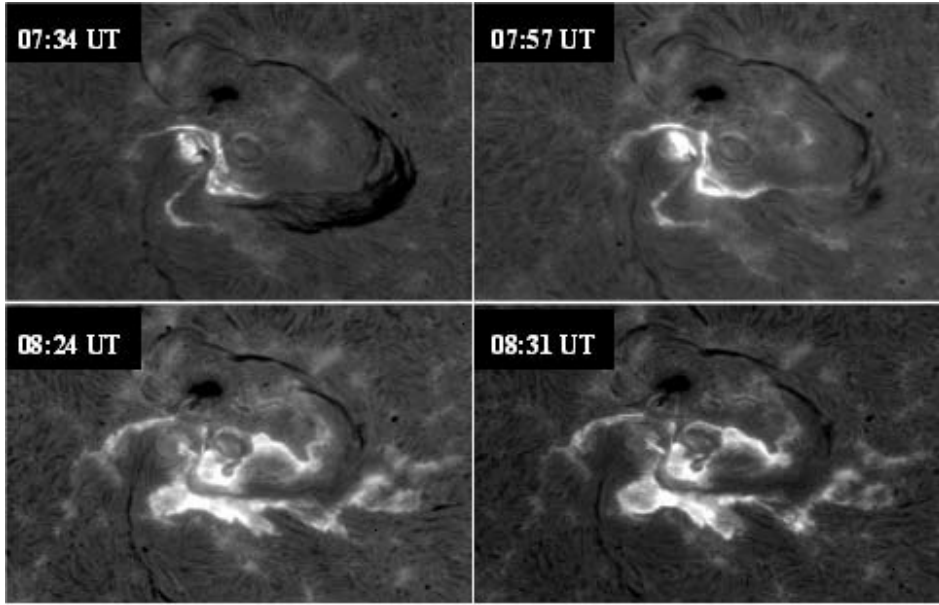


Figure 1. Time-lapse images of an M3.9, 2N flare recorded in H-alpha at Udaipur Solar Observatory. This flare was associated with the large CME of November 18, 2003.

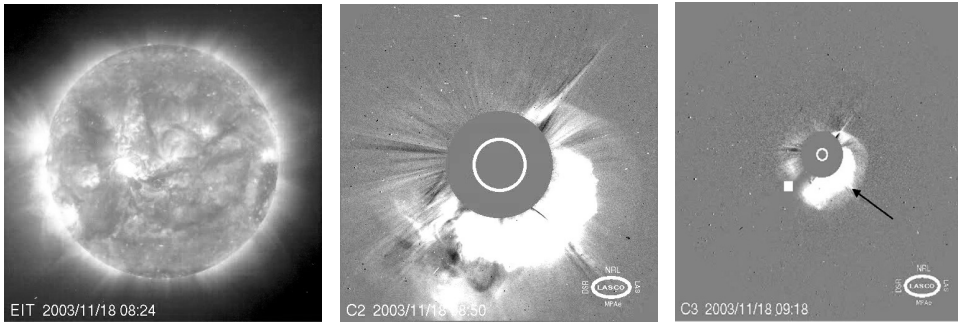


Figure 2. CME of November 18, 2003 which was the source of the strongest storm of the current solar cycle. The first image shows the flare as observed in EIT (195 Å) on SOHO. The middle and right images are recorded in white light by LASCO-C2 and C3.

shocks which were recorded by CELIAS aboard SOHO (<http://umtof.umd.edu/pm>). The corresponding IP shock velocities ranged between 730 and 2000 km s⁻¹. A deceleration of $\sim -3.3 \text{ m s}^{-2}$ was observed for the CME of November 18, 2003 as it travelled through the IP medium. On the other hand, the CME of October 28, 2003 is an example of faster CME with a relatively lower value of deceleration $\sim -1.8 \text{ m s}^{-2}$. The arrival time of November 18 CME was ~ 48 hours as compared to 19 hours in the case of October 28 CME, nevertheless, the magnitude of the resulting storm was greater in the case of the former making it the strongest storm of the current solar cycle ($D_{ST} \sim -472 \text{ nT}$). Although our previous studies indicate that the initial speed of a CME is one of the key factors that dictate the magnitude of the resulting geomagnetic

storm (Srivastava & Venkatakrishnan 2002, 2004), the November 18 event is an exception in that, it had a relatively low initial speed. However, the CMEs of October 28 and 29 had high initial speeds. This further suggests that there may be other solar factors that influence the magnitude of the resulting geomagnetic storms.

With exception of the November 18, 2003 CME, the initial speed of a CME may be considered as the most reliable predictor (amongst solar variables) of the strength of the associated geomagnetic storm because fast mass ejections are responsible for building up the ram pressure at the earth's magnetosphere. This result is based on a comparative study of these events with other super-storms of the current solar cycle (Srivastava 2005a). This study also showed the following:

- If the source active region of a CME is large in area, it possesses a large amount of magnetic energy. The larger the magnetic energy available, the higher the speed of the ensuing CME which causes higher ram pressure of the solar wind on the magnetosphere and compression of the southward component of the IMF (B_z).
- The strength of a geomagnetic storm not only depends on the interplanetary-magnetospheric coupling parameter VB_z , but also on the duration (T_{B_z}) of B_z . The higher the value of $VB_z T_{B_z}$ the higher will be the strength of the storm.

3. Space weather prediction: Development of logistic regression model

Solar and interplanetary parameters as described in the previous section were specified for 64 geo-effective CMEs observed during 1996–2002 and were used to develop a logistic regression model to predict the occurrence/non-occurrence of a severe storm. The geomagnetic storms with $-200 < D_{ST} < -100$ nT were classified as 'intense' and other events with $D_{ST} < -200$ nT as 'super-intense'. It may also be pointed out here, that the 'super-storms' with $D_{ST} < -300$ nT as observed during October–November 2003, are very rare. Therefore, the above classification was necessary, in order to have at least a sufficient number of events in both the categories so that a successful predictive model can be developed.

In the logistic regression model, a binary dependent variable indicating the occurrence of an intense or super-intense geomagnetic storms is regressed against a series of independent model variables that define a number of solar and interplanetary properties of the 64 geo-effective CMEs. The input variables for these models were chosen from Srivastava & Venkatakrishnan (2004). The value of the independent variable was taken as 1 for super-intense storms ($D_{ST} < -200$ nT) and 0 for intense storms ($-200 < D_{ST} < -100$ nT). The logistic regression equation is:

$$P = \frac{1}{(1 + e^{-Z})}, \quad (1)$$

where

$$Z = b_0 + b_1 x_{i1} + \cdots + b_j x_{ij},$$

P is the probability of the occurrence of an intense/super-intense geomagnetic storm, given the i th observation of the solar and interplanetary variables, and Z_i is the value of the continuous variable Z . The main solar inputs to the model included:

- latitude and longitude of the origin of the CME,
- flare/prominence association,

- association with full/partial/non-halo CME and
- initial speeds of the CME (V_i).

The interplanetary inputs were:

- shock speeds (V_{SH}),
- ram pressure (P_R),
- total value of the IMF (B_T),
- southward component of the IMF (B_z),
- solar wind speeds (V_1 and V_2) and densities (n_1 and n_2) before and after the shock and
- solar wind-magnetospheric coupling parameter (VB_z).

The details of the model have been described in Srivastava (2005b). The training of the model comprised estimation of regression coefficients using an iterative maximum likelihood method on a training sub-set of 55 geo-effective events. The trained logistic regression model was validated by predicting the occurrence of intense/super-intense geomagnetic storms for the validation sub-set of 9 events which were not used for training the model. The results show that the logistic regression model correctly classifies 62.5% of the training super-intense storms and 97% of the intense storms. Amongst the validation events, the model correctly classifies 50% of the super-intense storms and 100% of the intense storms.

4. Summary and conclusion

The major geomagnetic storms are mostly associated with CMEs accompanied by flares originating from large active regions located close to the central meridian and at low latitudes. Although the speed of the November 18, 2003 CME was relatively small, the resulting storm of November 20 possibly owes its large magnitude to the long duration for which the B_z component remained southward and to the large magnitude of B_z resulting from the high inclination of the magnetic cloud to the ecliptic plane. Further, if the source active region of a CME is large in area, it possesses a large amount of magnetic energy and hence is capable of producing a high speed CME. The high speed CME is responsible for causing higher ram pressure of the solar wind on the magnetosphere and compression of the southward component of the IMF (B_z). However, it does not imply that if the area is smaller then the ensuing CME cannot be strongly geo-effective, for example, the November 18, 2003 case shows that such CMEs may also lead to severe storms and arrive without prior warning. The strength of a geomagnetic storm not only depends on the interplanetary-magnetospheric coupling parameter VB_z but also on the duration of B_z . The higher the value of VB_z T_{B_z} the higher will be the strength of the geomagnetic storm (Srivastava & Venkatakrishnan 2004; Srivastava 2005a). Amongst the solar variables, the initial speed of a CME can be considered as the most reliable predictor of the strength of the associated geomagnetic storm. The initial speed not only influences the ram pressure but is also responsible for strengthening the interplanetary-magnetospheric coupling parameter VB_z .

Using these results as inputs, a simple logistic regression model has been developed. This model successfully predicts the occurrence of intense geomagnetic storms, although it is only moderately successful in predicting super-intense storms.

The results also indicate that as compared to the solar variables, the interplanetary variables are better predictors of the occurrence of geomagnetic storms, which is possibly due to the fact that the relevant solar variables for the prediction of the geomagnetic storms have not been identified as yet. Therefore, one needs to identify key solar variables that can be effectively used to predict the occurrence of a geomagnetic storm. A key solar variable may be the solar surface magnetic field, and its orientation, which can influence the interplanetary magnetic field and its orientation. Further, some of properties of CMEs may get modified due to interaction with the ambient medium as they propagate through the interplanetary medium, which may also affect the predictive capability of the model. These factors will be taken into account for improving the predictive model, in future.

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Coronal Mass Ejections of Solar Cycle 23

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Abstract. I summarize the statistical, physical, and morphological properties of coronal mass ejections (CMEs) of solar cycle 23, as observed by the Solar and Heliospheric Observatory (SOHO) mission. The SOHO data is by far the most extensive data, which made it possible to fully establish the properties of CMEs as a phenomenon of utmost importance to Sun–Earth connection as well as to the heliosphere. I also discuss various subsets of CMEs that are of primary importance for their impact on Earth.

Key words. Sun: coronal mass ejections, shocks, magnetic field.

1. Introduction

Coronal mass ejections (CMEs), as we know them today, were first detected in the coronagraph images obtained on December 14, 1971 by NASA's OSO-7 spacecraft (Tousey 1973). Typical coronagraphs have an occulting disk to artificially eclipse the bright photosphere, so the faint coronal structures outside the periphery of the occulting disk can be viewed in the photospheric light scattered off these structures. CMEs are now understood as large-scale magnetized plasma structures originating from closed magnetic field regions on the Sun: active regions, filament regions, active region complexes and trans-equatorial interconnecting regions. Definite inferences on mass ejections in the solar atmosphere predated the white-light discovery by decades: prominence eruptions (Secchi and de la Rue in the late 1800s (see, e.g., Tandberg-Hanssen 1995)), slow-drifting radio bursts (Payne-Scott *et al.* 1947), and moving type IV radio bursts (Boischot 1957). There were also other indications around the time of the white light detection: rapid decay of transient coronal condensation (Hansen *et al.* 1971), coronal green line transients (DeMastus *et al.* 1973), helium abundance enhancements in the high speed plasmas behind interplanetary (IP) shocks (Hirshberg *et al.* 1972) later recognized as the IP counterpart of CMEs (Borrini *et al.* 1982). CMEs also existed conceptually as agents causing geomagnetic storms (Lindemann 1919) and driving IP shocks (Gold 1955). A list of review articles covering the three decades of CME research can be found in Gopalswamy (2004). Here, I highlight the recent research on CMEs spurred by the extensive and uniform data from the Solar and Heliospheric Observatory (SOHO) mission's Large Angle and Spectrometric Coronagraph (LASCO, Brueckner *et al.* 1995).

2. CME properties

The basic attributes of a CME are its speed, width, acceleration, and central position angle (CPA), all with reference to the sky plane. These are obtained from a time sequence of coronagraphic images, in which the CME can be recognized as a moving feature occupying a well-defined region. The angular extent of the moving feature defines the width. The central angle of this extent with reference to the solar north is the CPA. The speed is normally determined from a linear fit to the height–time (h – t) plots. But CMEs often have finite acceleration, so the linear-fit speed should be understood as the average value within the coronagraphic field of view. Quadratic fit to the h – t plots gives the constant acceleration, which again is an approximation because the acceleration may also change with time.

The measured sky-plane speed ranges from a few km s^{-1} to $\sim 3000 \text{ km s}^{-1}$ (see e.g., Gopalswamy 2004; Yashiro *et al.* 2004), with an average value of $\sim 483 \text{ km s}^{-1}$. The CME speed has a lognormal distribution (Yurchyshyn *et al.* 2005). The apparent angular width of CMEs ranges from a few degrees to more than 120 degrees, with an average value of $\sim 46 \text{ deg}$. The average width was computed for CMEs with width $\leq 120 \text{ deg}$. The true width of halo CMEs (Howard *et al.* 1982), which appear to surround the occulting disk, is unknown. Most of the h – t plots fall into three types: accelerating, constant speed, and decelerating, indicating different degrees of propelling and retarding forces acting on CMEs (Gopalswamy *et al.* 2001a). Figure 1 shows a scatter plot between the measured acceleration (a in ms^{-2}) and speed (v in km s^{-1}) of all the CMEs for which the acceleration estimate was possible. Despite the large scatter, the acceleration has a reasonable correlation with speed: $a = -0.0133(v - 452)$, which shows that slow CMEs ($v < 450 \text{ km s}^{-1}$) accelerate, CMEs of intermediate speed ($v \sim 450 \text{ km s}^{-1}$) have no appreciable acceleration, and fast CMEs ($v > 450 \text{ km s}^{-1}$) decelerate. Close to the Sun, a is determined by the propelling force, gravity and coronal drag. Note that $a = 0$ for $v = 452 \text{ km s}^{-1}$ means that the solar wind has already picked up speed within the LASCO field of view. The CME acceleration over the entire Sun–Earth distance also has a similar relation to the CME initial speed (Gopalswamy *et al.* 2001b): $a = -0.0054(v - 406)$. In this case, the propelling force and gravity become less important and the drag force dominates so that $v = 406 \text{ km s}^{-1}$ corresponds to the average solar wind speed.

The width is a good indicator of the mass content of CMEs. The mass is estimated as the excess mass in the coronagraphic field of view assuming that the entire mass is located in the sky plane (see, e.g., Vourlidas *et al.* 2002). Like the width, the mass also changes during the early phase of the CME before stabilizing to a near-constant value. This constant value is used as the representative mass. Figure 2 shows that the mass ranges from a few times 10^{13} g to more than 10^{16} g . The kinetic energy obtained from the measured speed and mass ranges from $\sim 10^{27} \text{ erg}$ to $\sim 10^{32} \text{ erg}$, with an average value of $5 \times 10^{29} \text{ erg}$. Some very fast and wide CMEs have kinetic energies exceeding 10^{33} erg , generally originating from large active regions [Gopalswamy *et al.* 2005a]. Most of these results are similar to those obtained by pre-SOHO coronagraphs, except for the smaller average mass and a larger number of fast and wide CMEs, including halo CMEs.

CMEs often display spatial structures, commonly referred to as the ‘three-part structure’ (a frontal structure, central bright core, and cavity – see Hundhausen 1999). The frontal structure overlies the cavity, which contains the bright core. The core has been

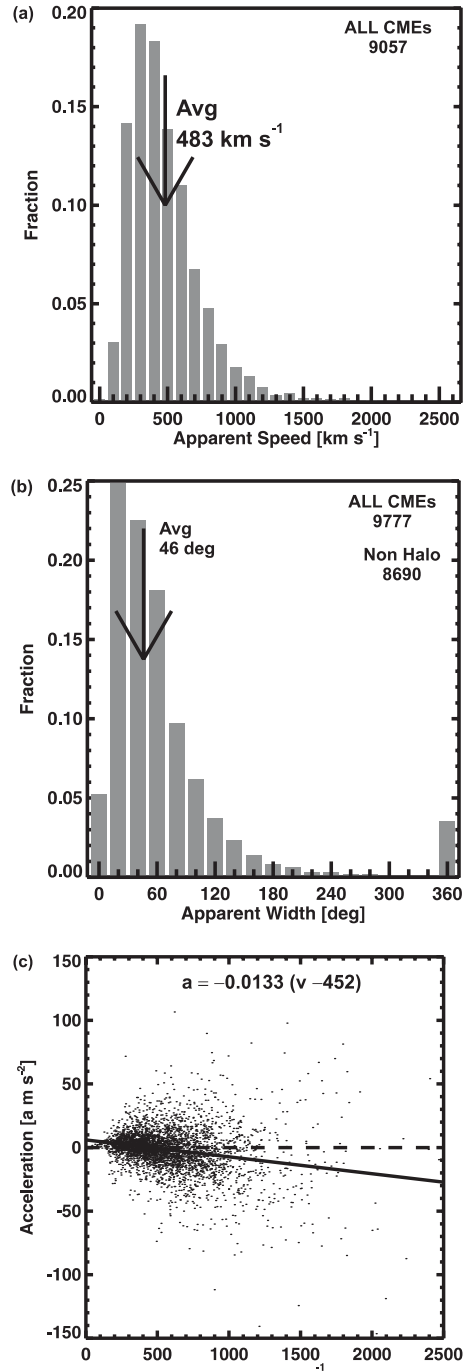


Figure 1. CME speed (a), width (b) and acceleration (c) of all CMEs from 1996 to June 2005. The acceleration has a large scatter, but there is a clear trend that the fast CMEs decelerate, while slow CMEs accelerate.

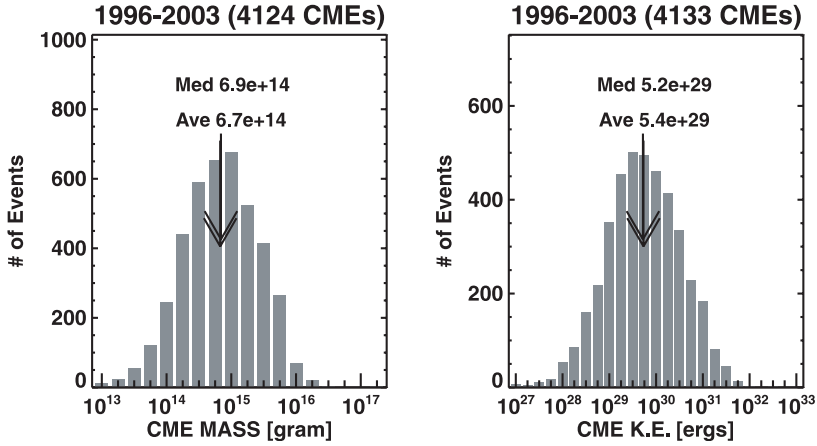


Figure 2. CME mass (left), and kinetic energy (right) of SOHO CMEs for the period 1996–2003.

shown to be the eruptive prominence by comparing coronagraph and H-alpha observations. Eclipse pictures often show the three-part structure in the pre-eruptive stage, where the helmet streamer takes the place of the frontal structure. Not all CMEs show the three-part structure either due to geometrical reasons (Cremades *et al.* 2004) or due to the nature of the source region. Halo CMEs often contain a diffuse front with a bright inner feature of smaller spatial extent. The cavity, if any, must have been obscured by the overlying structure. If the CME is fast with respect to the local Alfvén speed, it drives a shock. Such CMEs must have four-part structure, but the shock is often difficult to detect. The circular shape of the outermost structure in halo CMEs may indeed be a manifestation of the shock (Sheeley *et al.* 1999).

The CME plasma is multithermal with the prominence core at ~ 8000 K and the outer structure at a few MK. Occasionally, CMEs may consist of flare ejecta with temperature exceeding 10 MK. The magnetic field in CMEs is not directly measurable near the Sun. The magnetic field in prominences is typically up to 30 G, while it can exceed 1000 G in the active region cores. The field strength in the outer corona is typically less than ~ 1 G. The coronal cavity overlying the filament may also have field strengths higher than the overlying corona for pressure balance requirements. The 1-AU field strength in CMEs is typically tens of nT, which is up to an order of magnitude higher than the field in the quiet solar wind (a few nT). Tens of nT at 1 AU correspond to tens of G at the Sun if we assume a simple r^{-2} dependence. This is consistent with the field strength in the prominence and the inferred values in the overlying coronal cavity.

3. CME rate and implications

How frequently do CMEs occur? During solar minima, one CME occurs every other day. The rate goes up to several per day during solar maximum. On one day during solar maximum, 13 CMEs were recorded by SOHO; there were several days with more than 10 CMEs (Gopalswamy *et al.* 2003a). The daily CME rate averaged over Carrington Rotations (27.3 days) was found to exceed 6/day (see Fig. 3). During cycle 23, the rate

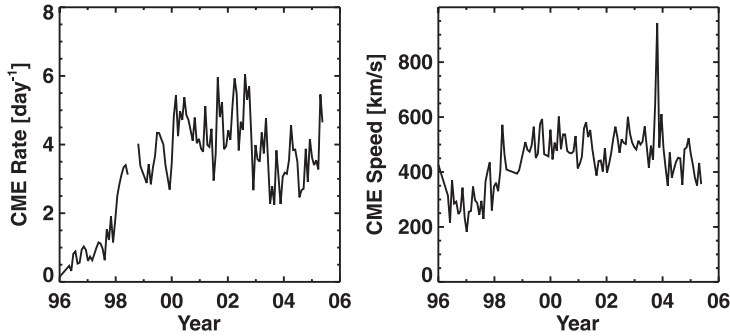


Figure 3. CME rate (left) and mean speed (right) averaged over Carrington Rotation (CR) periods. The gap in the CME rate is due to the temporary disability of SOHO during June–October 1998. The large spike in the CME mean speed is due to the Halloween 2003 CMEs.

increased abruptly in 1998 and remained relatively high through the middle of 2002. There is a general decline after 2002, but the rate is still much higher than the solar minimum values (as of mid 2005). The large spikes in Fig. 3 are due to superactive regions, which are prolific producers of CMEs.

The pre-SOHO CME rate was found to have a good correlation with the sunspot number (SSN), which indicated a solar maximum rate of ~ 3 (Hildner *et al.* 1976), confirmed by data from the SMM and Solwind coronagraphs (Cliver *et al.* 1994). The SOHO CME rate, although initially consistent with pre-SOHO results (St Cyr *et al.* 2000), has several unexpected aspects:

- The solar maximum rate nearly doubled to 6/day.
- While SOHO data confirmed the SSN-CME rate relationship, the correlation is less than perfect, which can be attributed to the high-latitude (HL) CMEs associated with the polar crown filaments (PCFs) around solar maximum. The HL CMEs are not associated with sunspots (Gopalswamy *et al.* 2003a,b) and hence the weaker correlation.
- The SSN and CME rate reached maximum at different peaks of the double maximum (late 2000 for SSN and mid 2002 for CME rate). This result also follows from the fact that the HL CMEs are associated with PCF (spotless eruptions), while the low-latitude (LL) CMEs are predominantly from sunspot regions (except for some CMEs associated with low-latitude quiescent prominences).
- There is a remarkable coincidence between the cessation of HL CMEs and the polarity reversal at solar poles. This result holds good for individual poles, which reversed at different times (north in July 2000 and south in May 2002) (Gopalswamy *et al.* 2003b).
- The rate of CMEs and the minimum-to-maximum variability, originally thought to be inadequate (Newkirk *et al.* 1981; Wagner 1984), have been found to be large enough to cause galactic cosmic ray modulation (Gopalswamy 2004; Lara *et al.* 2005). HL CMEs have been found to be more effective in suppressing the cosmic rays during the $A > 0$ epoch (when solar north pole has north polarity), while the LL CMEs seem to be dominant in the $A < 0$ epoch (when solar north pole has south polarity). The galactic cosmic rays enter the heliosphere from the polar

regions during the $A > 0$ epochs (encountering HL CMEs on their path), while they enter in the equatorial region during the $A < 0$ epochs (encountering the LL CMEs).

Another CME attribute linked with solar activity is the mean speed, which doubled from minimum to maximum. This was not firmly established during cycle 22 (Webb & Howard 1994; Hundhausen 1999; Gopalswamy *et al.* 2003a). The mean speed also shows the double peak but the difference between the first and second peaks is not significant. The huge spike in the mean speed is due to the famous Halloween events (October–November 2003) when an extraordinary set of fast CMEs occurred from three active regions (AR 10484, 10486, and 10488) that had significant consequences throughout the heliosphere (see Gopalswamy *et al.* 2005b for a list of publications on these events).

4. CMEs and space weather

CMEs are large-scale magnetized plasma structures, so their propagation into the heliosphere has important consequences. Earth-directed CMEs are likely to impact the magnetosphere to cause geomagnetic storms. CME-driven shocks accelerate solar energetic particles (SEPs). Gosling (1993) had pointed out the central role played by CMEs in causing intense geomagnetic storms. This became further clear when the direct connection between CMEs and their interplanetary counterparts (ICMEs) was established (Burlaga 1981; Lindsay *et al.* 1999; Gopalswamy *et al.* 2000). Halo CMEs, once thought to be of questionable existence, have proved to be the major players in space weather, thanks to the high sensitivity and large field of view of the SOHO/LASCO coronagraphs (see, e.g., Webb *et al.*, 2000). The geoimpact of CMEs generally falls into two categories: geoeffectiveness and SEPeffectiveness. Geoeffective CMEs cause non-recurrent (also known as transient) geomagnetic storms. SEP-effective CMEs cause the gradual and long-lasting SEP events with an intensity of at least 10 pfu in the >10 MeV channels of particle detectors such as on GOES.

4.1 Geoeffective CMEs

CMEs originating from close to the disk center (most of which become front-side halos) directly impact Earth and produce geomagnetic storms provided their magnetic field has a southward component (B_s). The impact is marked by the sudden commencement if the CME drives a shock. If the shock sheath region has a B_s , the storm starts immediately after the shock. ICMEs with a flux rope structure (magnetic cloud), almost always cause a storm because either the front or the rear section of the ICME contains B_s . Occasionally CMEs arrive at Earth with a high inclination resulting in an intense storm (when the magnetic field is fully southward, see Gopalswamy *et al.* 2005c) or no storm at all (when the magnetic field is fully northward, see Yurchyshyn *et al.* 2001). The CME speed and the strength of the magnetic field it contains primarily decide the intensity of the geomagnetic storms: Figure 4 shows that the product of CME speed (V_{cme}) and the 1-AU field strength (B) has the best correlation with Dst index. Various combinations V_{cme} , ICME speed (V_{mc}), B , and B_s have been considered in the past (Yurchyshyn *et al.* 2004; Srivastava & Venkatakrishnan 2004 and references therein), but most of the correlation coefficients are generally lower than the one given

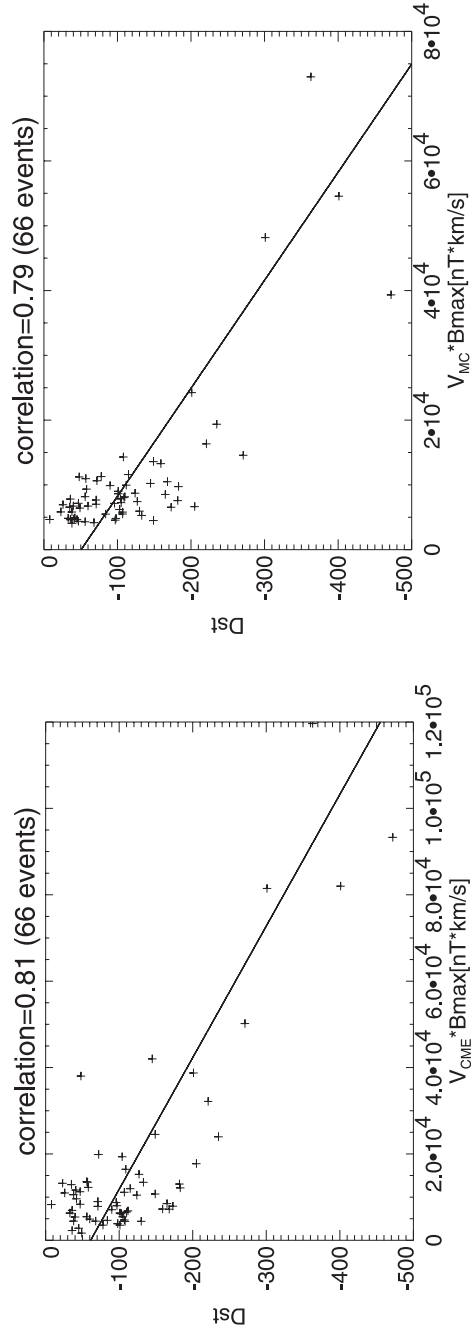


Figure 4. Correlation $V_{CME}B$ (left) and $V_{MC}B$ (right) with Dst (nT), where B is the 1-AU magnetic field in CMEs, V_{CME} is the CME speed near the Sun and V_{MC} is the speed of the magnetic clouds at 1 AU. The suffix “max” indicates the peak value measured at 1 AU.

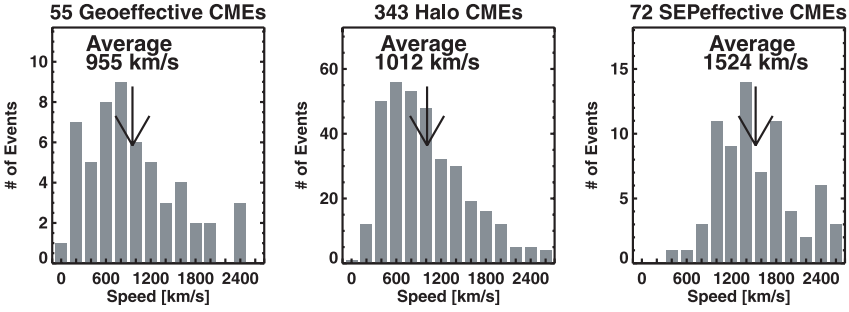


Figure 5. Speed distributions of geoeffective CMEs (left), Halo CMEs (middle) and SEPeffective CMEs (right). The mean speeds are marked.

in Fig. 4. Of course, the best set of parameters will be V_{cme} and the CME magnetic field (B_{cme}) measured near the Sun. We are far from obtaining B_{cme} , so we used the 1-AU B as proxy.

4.2 SEPeffective CMEs

The gradual SEP events are thought to originate from CME-driven shocks (Reames 1999; Kahler 2001; Gopalswamy *et al.* 2003c; 2004). The associated CMEs are fast and wide (average speed $\sim 1500 \text{ km s}^{-1}$, mostly full or partial halos – see Fig. 5). SEP events with ground level enhancements (GLEs) are associated with the fastest of all CMEs (average speed $\sim 2000 \text{ km s}^{-1}$, see Gopalswamy *et al.* 2005d). Up to 15% of the CME kinetic energy goes into the accelerated particles (Emslie *et al.* 2004), suggesting that the CME-driven shocks are efficient particle accelerators. The SEP events are also closely related to type II radio bursts (Gopalswamy 2003; Cliver *et al.* 2004) since the same shocks accelerate SEPs and electrons. There is also a hierarchical relationship between type II bursts and CME kinetic energy (Gopalswamy *et al.* 2005e): CMEs with above-average kinetic energy produce type II bursts confined to the metric (m) domain; CMEs with moderately high kinetic energy produce type II bursts in the decameter-hectometric (DH) domain; CMEs with the highest kinetic energy produce type II bursts over all the wavelength domains (metric to kilometric or mkm). There is also a small population of purely km type II bursts due to accelerating CMEs that drive shocks only at large distances from the Sun. It is not surprising that most of the CMEs associated with mkm type II bursts are also associated with SEP events (see also Fig. 6).

4.3 Comparing geoeffective and SEPeffective CMEs

Figure 5 compares the speeds of geoeffective CMEs, halo CMEs, and SEPeffective CMEs. Clearly the halo and geoeffective CMEs have the same average speed, as expected. The SEPeffective CMEs, on the other hand, are much faster. One might wonder why the number of geoeffective CMEs is only $\sim 1/6$ of the number of halos. First of all, only $\sim$ half of the halos are front-sided. Only about a third of the front-sided halos are likely to originate within 30° of the disk center and this number is

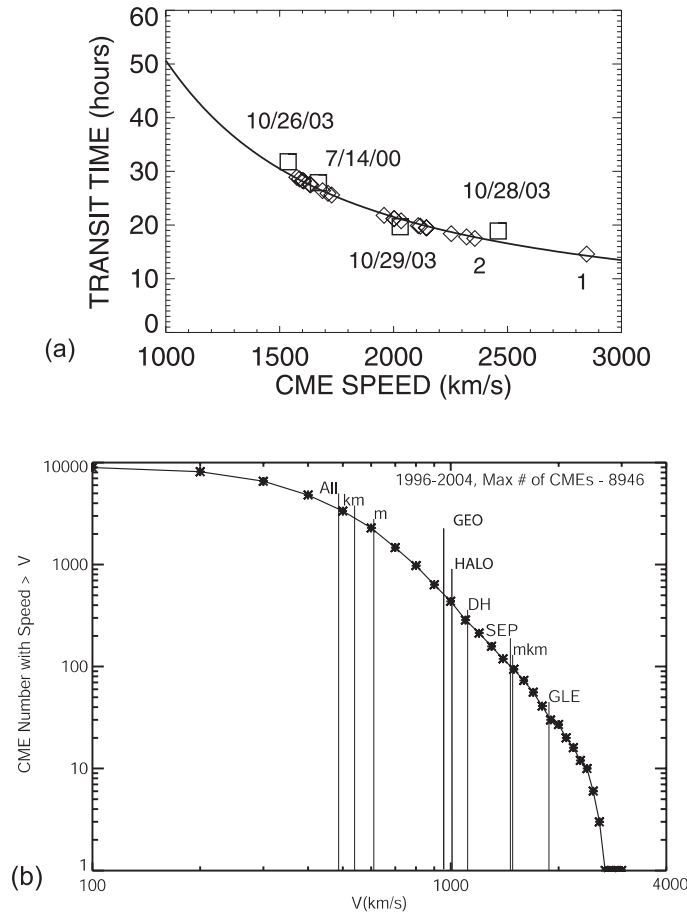


Figure 6. (a) The Sun–Earth transit time of CME-driven shocks according to the ESA model approximates as $T = ab^V + c$, where $a = 151.002$, $b = 0.998625$, and $c = 11.5981$ (solid curve). The open squares are fast-transit events of cycle 23. The diamonds are the historical events, with the two fastest events marked: (1) August 4, 1972 and (2) September 1, 1859. (b) The number of CMEs above a certain speed, V (cumulative lognormal distribution) with the average speeds of various CME populations marked.

similar to the number of geoeffective CMEs. Most of the false alarms are caused by asymmetric halos (which originate beyond 30 deg from disk center). Another cause may be merging between successive CMEs. The source regions of the SEP-effective CMEs are generally located on the western hemisphere, although occasionally they do originate from the eastern hemisphere (Gopalswamy *et al.* 2005). The geoeffective CMEs, on the other hand, originate close to the disk center. Thus, the union of SEP- and geoeffective CMEs (most of the front-sided fast and wide CMEs) is important in deciding the conditions in Earth's space environment; their intersection defines the set of CMEs having multiple effects (producing both SEP events and geomagnetic storms).

4.4 Sun–Earth transit of geoeffective CMEs

The speed distribution of the geoeffective CMEs gives an indication of the time taken by the CME-driven shocks to arrive at Earth. Figure 6 (left) shows the travel time of shocks driven by CMEs with speeds $\geq 1000 \text{ km s}^{-1}$. The solid curve is the empirical shock arrival (ESA) model, which relates the Sun–Earth transit time (T in hours) to the CME speed ($V \text{ km s}^{-1}$) by the formula, $T = ab^V + c$, where $a = 151.002$, $b = 0.998625$, and $c = 11.5981$ (Gopalswamy *et al.* 2005a). T ranges from 50 h for 1000 km s^{-1} CMEs to 13.5 h for 3000 km s^{-1} CMEs. Also shown are some fast-transit shocks from SOHO (the Bastille Day shock of 2000 and three shocks from the Halloween 2003 period). The transit times of historical fast-transit events, whose CME speeds are inferred from the above formula, are also shown for comparison. The transit-time formula suggests that the Sun–Earth transit times of shocks may not exceed ~ 12 h, because the number of CMEs with speeds exceeding 2500 km s^{-1} drops precipitously (see the cumulative distribution shown in Fig. 6).

In summary, we need to worry only about a subset of CMEs for space weather purposes:

- CMEs associated with DH and mkm type II bursts,
- SEP-producing CMEs,
- geoeffective (or halo) CMEs. Figure 6 (right) shows that there are only ~ 1000 CMEs ($\sim 10\%$ of all CMEs) that fall into this group. CMEs associated with SEPs having ground level enhancement (GLE) have the largest speed, while the CMEs associated with purely metric (m) and purely kilometric (km) type II bursts have just above average speeds.

5. Concluding remarks

The high quality data over an extended period of time from SOHO have provided new insights into the problem of CMEs. We now recognize that CMEs constitute the most energetic phenomenon in the heliosphere. While we have a good idea on the kinematic properties of CMEs, their magnetic properties, especially close to Sun, are poorly understood. It is ironic that the magnetic properties are primary indicators of CMEs at 1 AU, but near the Sun, there is very limited information. Understanding the vector magnetic fields of CMEs near the Sun, and their relation to surface and subsurface evolution is crucial in developing long-term prediction of CME production. This will also aid the prediction of potential geoimpact because the magnetic field orientation is a critical parameter. We have seen that only the fast and wide CMEs ($\sim 10\%$ of all the CMEs) have space weather implications. Therefore, modeling efforts should concentrate on these CMEs for space weather applications. SOHO has provided a great deal of information on the initiation and 3D structure of CMEs. Yet, connecting disk signatures to coronagraphic signatures is still not easy due to the occulting disk employed by the coronagraphs. The STEREO mission is likely to alleviate some of these difficulties. STEREO has also the capability to provide information on the early-phase acceleration of CMEs. While the connection between SEP events and CMEs is fairly clear, there are many fast and wide CMEs that are not associated with SEPs. This needs a better characterization of the ambient medium such as density and magnetic field, which determine the local Alfvén speed and hence shock formation. The effects

of CME-induced turbulence and successive CMEs in the ambient medium have to be incorporated into the shock-acceleration theories.

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Photospheric, Chromospheric and Helioseismic Signatures of a Large Flare in Super-active Region NOAA 10486

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Abstract. NOAA 10486 produced several powerful flares, including the 4B/X17.2 superflare of October 28, 2003/11:10 UT. This flare was extensively covered by the H_α and GONG instruments operated at the Udaipur Solar Observatory (USO). The central location of the active region on October 28, 2003 was well-suited for the ring diagram analysis to obtain the 3-D power spectra and search for helioseismic response of this large flare on the amplitude, frequency and width of the p-modes. Further, using USO observations, we have identified the sites of new flux emergences, large proper motions and line-of-sight velocity flows in the active region and their relationship with the flare.

Key words. Solar flares—flare trigger—ring diagram.

1. Introduction

Solar flares are usually associated with prominent chromospheric and coronal signatures, and in exceptional events even at the photospheric level. The basic flare eruption is believed to occur due to a magnetohydrodynamic (MHD) catastrophe, driving a reconnection process in the magnetic field lines stretched out by the eruption (Priest & Forbes 2002). Extensive efforts have been made to identify the conditions and magnetic topology leading to the onset of flares and CMEs, viz., sigmoid formation along quasi-separatrix layers (Fletcher *et al.* 2001; Gibson *et al.* 2002), magnetic helicity injection by photospheric motions (Zhang & Wang 2002a), rapid magnetic field changes before the onset of flare (Wang *et al.* 2002; Spirock *et al.* 2002), magnetic shearing (Ambastha *et al.* 1993), intrusion of adjacent magnetic feature (Ulrich *et al.* 2002), magnetic flux emergence (Zhang & Wang 2002b; Kurokawa *et al.* 2002), etc. Also, moderate flares may occur where newly emerged flux gets cancelled with opposite polarity, presumably by reconnection (Martin *et al.* 1984; Ambastha & Mathew 2000; Jennings *et al.* 2002), and can trigger the release of magnetic energy stored in a larger volume (Priest 1984).

During October 18–November 5, 2003, three large sunspot groups, NOAA 10484, 10486 and 10488, were observed which produced 140 flares; 11 of which were X-class. Super-active region NOAA 10486 accounted for most of these extremely energetic flares. Figure 1 shows the GOES soft X-ray flux profile for October 26–31, 2003 showing two of its major flares which occurred when NOAA 10486 was located close to the solar disk-center. The record-setting flare of November 4, 2003 occurred when

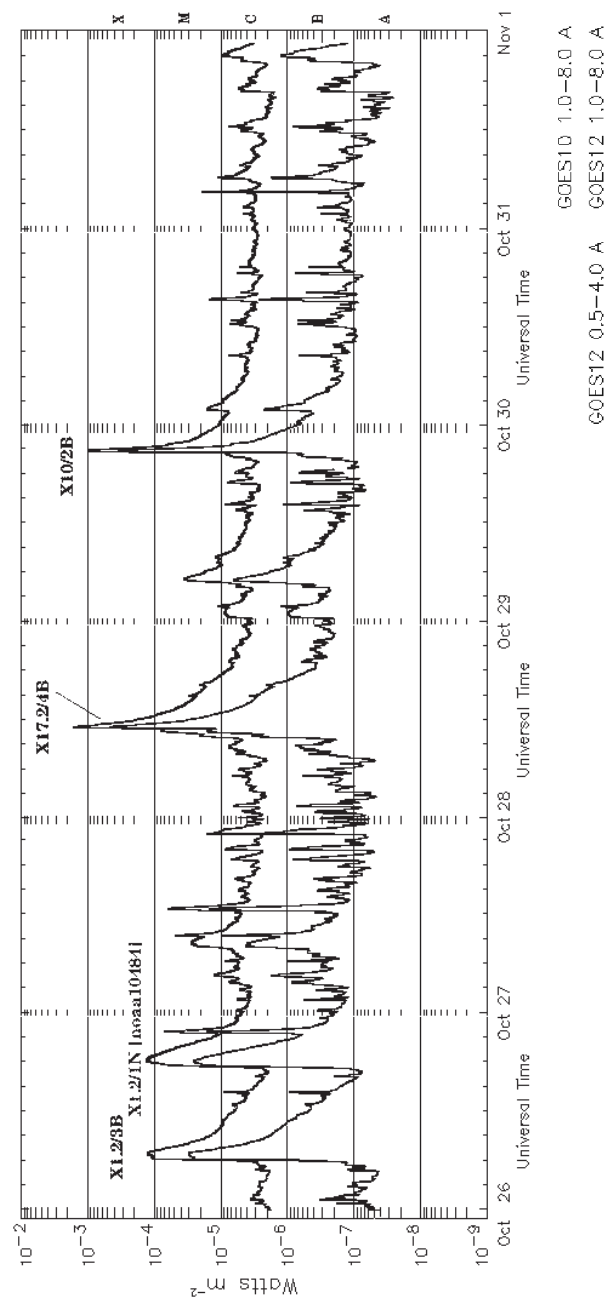


Figure 1. GOES disk-integrated soft X-ray flux showing the two major flares, X17.2 and X10, of NOAA 10486 which occurred during October 26–31, 2003.

NOAA 10486 reached the west-limb. This flare of unprecedented magnitude which saturated the GOES X-ray detectors was first classified as X28/3B flare, but later re-classified as an $X45 \pm 5$ flare based on its ionospheric response (Thomson *et al.* 2004). Most of these energetic flares were associated with fast moving CMEs, which continued to occur even as the active region rotated off the west-limb. Due to its central location on the disk, the 4B/X17.2 flare event of October 28, 2003/11:10 UT and associated proton storm and CME had much larger geomagnetic effects as compared to the larger X45 flare which occurred at the limb. The total solar irradiance (TSI) due to this flare recorded an increase by 360 mW/m^2 , the first such unambiguous detection since the beginning of routine TSI measurements in 1978 (Woods *et al.* 2004). We have selected this flare for detailed analysis of its photospheric, chromospheric and helioseismic signatures due to the favourable location of the active region near the CMP which minimizes the projection effects.

2. The observational data

High cadence H_α filtergrams were obtained from USO providing extensive coverage of the active regions NOAA 10484, 10486 and 10488 during the period of October 18–November 5, 2003. In addition, magnetograms and dopplergrams were also obtained at a cadence of 1-minute by the GONG+ instrument operated at USO. Using movies made from these observations, we attempt to identify the sites of magnetic flux changes, chromospheric restructuring, and the doppler velocity changes in the active region, particularly around the site of flare onset. The two other large regions, NOAA 10484 and 10488, were also observed during this period, however, they showed considerably lower flare productivity as compared to NOAA 10486.

3. Evolution of NOAA 10486 and the 4B/X17.2 flare

Figure 2 shows daily evolution of chromospheric structures of NOAA 10486 using H_α -filtergrams obtained from USO during October 26–31, 2003. A long curved filament ‘a’ was seen to meander through the middle of the active region on October 26, 2003 from which a smaller filament ‘b’ branched out. This system of filaments went through a major structural change on October 27, 2003 forming a ‘channel’ of opposite polarity region between them, marked as ‘c’. Yang *et al.* (2004) reported strong shear flows along the magnetic neutral line, i.e., the filament ‘c’. Zirin & Wang (1993) earlier discovered such magnetic ‘channel’ structures of elongated mixed polarity regions with strong shear flows along them. Harvey & Harvey (1976) have also found strong horizontal shear flows in flaring regions, and suggested them to be an important factor in flare production. This appears to be a common feature of all superactive regions. As seen from the filtergram of October 28, 2003, the channel gradually became narrower as a result of significant spatial changes taking place in the active region. Interestingly, this channel was the site of the onset of the 4B/X17.2 superflare of October 28, 2003 (Fig. 3). Before the 4B flare, several smaller flares occurred in the neighbouring locations. Chromospheric structural changes were observed subsequent to the superflare, for example, the filament ‘d’ disappeared, disconnection of ‘b’ occurred as marked at ‘e’ (Fig. 2: October 29 frame). New active filaments formed later during October 30–31, 2003 at the location ‘c’ indicating significant magnetic restructuring. The active

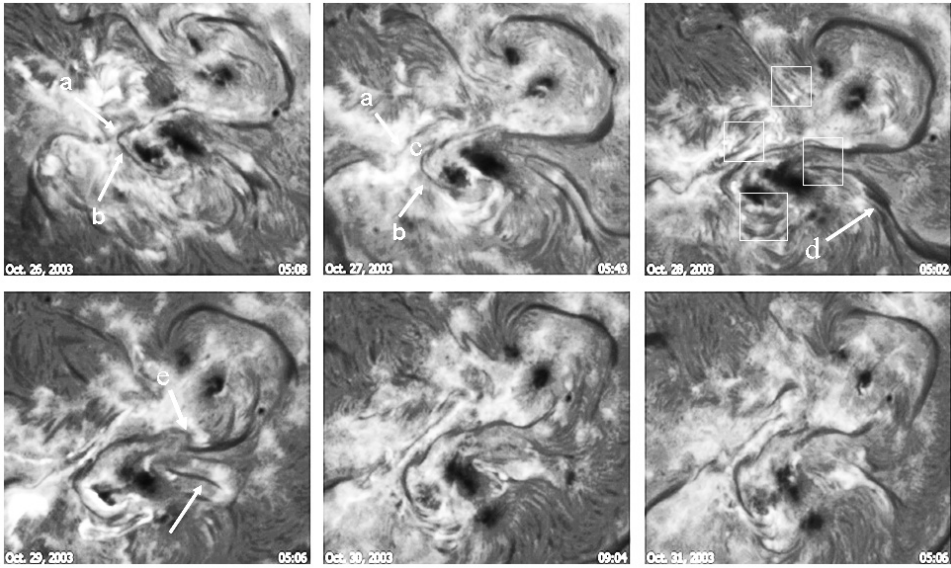


Figure 2. USO H_{α} filtergrams of NOAA 10486 during 26–31 October 2003 showing the evolution of chromospheric structures.

region retained its large magnetic complexity, and contained sufficient energy even after the superflare of October 28, 2003; evident from the fact that a large 2B/X10 white light flare followed at 20:37 UT on October 29, 2003. NOAA 10486 was estimated to possess an order of magnitude greater free energy than that in two other active regions NOAA 10484 and NOAA 10488 on October 29, 2003 (Metcalf *et al.* 2005).

Figure 3 shows the spatial and temporal evolution of the 4B/X17.2 super-flare in H_{α} . Negative filtergrams have been used for showing the structural details of the flare-ribbons more clearly. After the preflare activation seen in 10:39:28 UT frame, the flare erupted impulsively, reaching a maximum phase at 11:10:05 UT. The bright flare-ribbons separated away rapidly; a typical property of two-ribbon flares. Post-flare loops, generally observed during the decay phase, appeared as seen in the 11:14:46 UT frame, while the flare lasted over much longer. Notably, the H_{α} filtergrams showed formation of a twisted rope-like structure in the magnetic channel, rising upwards and eventual eruption of the filament.

4. *GONG+* magnetic and velocity field observations

Movies of *GONG+* velocity and line-of-sight magnetic field maps obtained during October 28, 2003, showed significant evolutionary changes at several locations along the magnetic channel and in the vicinity of the elongated filament. This is as shown in Fig. 4 in which L-O-S magnetogram contours, at ± 10 , ± 500 , ± 1500 G levels, are overlaid on the *GONG+* dopplergrams. Development of an increasing downward flow was seen in box ‘3’, where the filament was destabilized. On the other hand, enhancement of upward directed flow was observed in box ‘2’, where a positive polarity flux emerged. In box ‘1’, which covers location of the H_{α} twisted rope structure, moving blue shift events were observed. These evolved considerably during the time

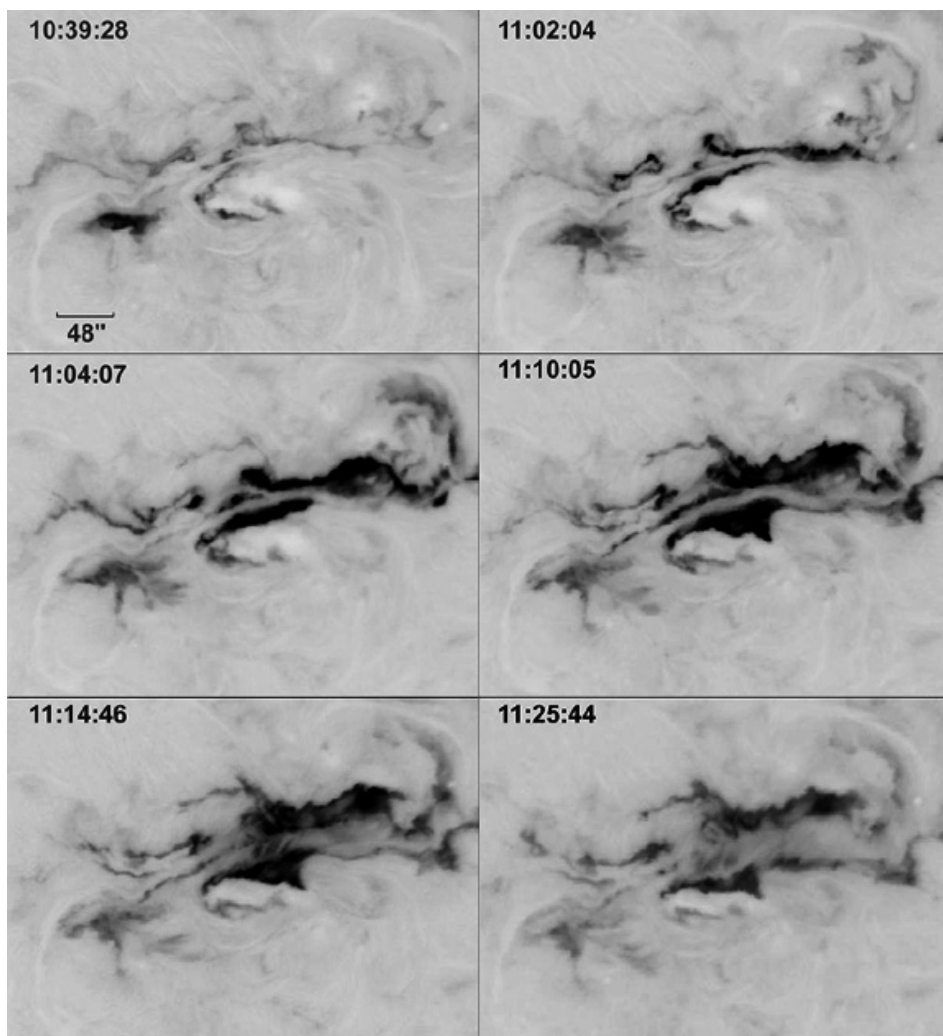


Figure 3. USO H_{α} filtergrams the 4B/X17.2 flare of October 28, 2003. Filtergrams are shown in 'negative'.

of the flare onset. Such events are indicators of upflow from reconnection events that are found to be 5 times more frequent before eruptive flares than in noneruptive flares (DesJardins & Canfield 2003), and upflows in the range of $40\text{--}80\text{ km s}^{-1}$ have been detected during the pre-flare phase (Brosius & Phillips 2004). Keil *et al.* (1994) found that flare kernels are locations of shear in vertical photospheric flows and locations of convergence in horizontal photospheric flows. Meunier & Kosovichev (2003) have also found evidence of persistent, supersonic downflows and shear flows in flaring active regions.

Several attempts have been made in the past to monitor the changes in magnetic flux, shear and electric currents before, during and after flares (Ambastha *et al.* 1993; Chen *et al.* 1994; Kosovichev & Zharkova 2001; Meunier & Kosovichev 2003; Yurchyshyn *et al.* 2004). Spirock *et al.* (2002) carried out a magnetic flux analysis for the large X20

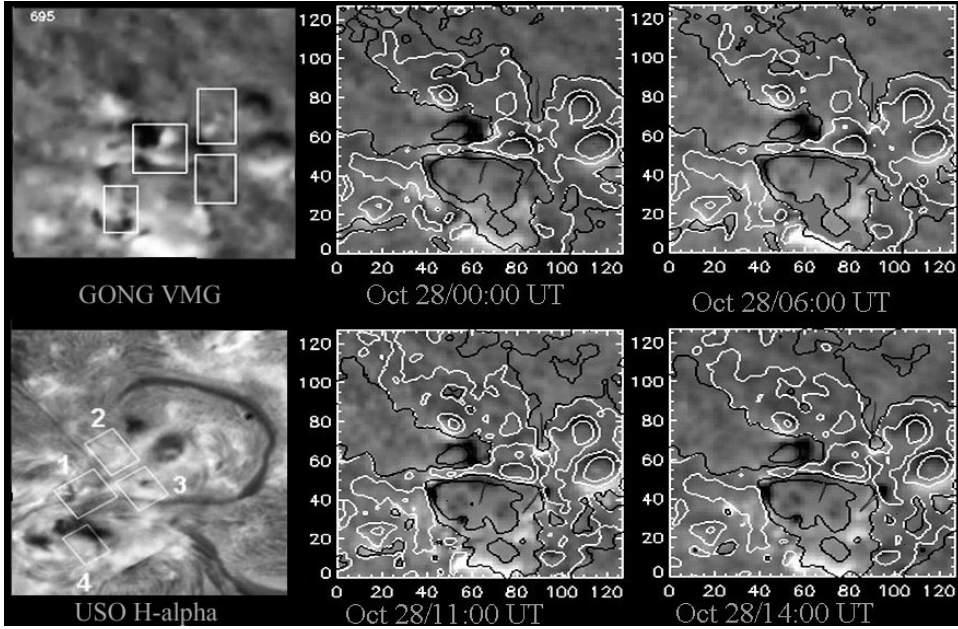


Figure 4. Overlays of GONG+ Dopplergrams (grey-level images) and line-of-sight magnetograms (contours). Selected areas-of-interest are shown over GONG VMG and USO H $_{\alpha}$ images.

flare of April 2, 2001, and reported an increase by 6×10^{20} Mx in the leading polarity, and no clear change in the following polarity. Wang *et al.* (2002) showed similar flux variation for six other X-class flares. However, the changes in the magnetic field twist parameter was statistically found insufficient to distinguish between flaring and non-flaring active regions (Leka & Barnes 2003). Recently, using GONG+ magnetograms, Sudol & Harvey (2005) have reported statistically significant changes in magnetic flux related with flares. However, it is important to consider flare-related changes in photospheric spectral lines while deriving such inferences, as this may seriously affect magnetic field calculations (Abramenko & Baranovsky 2004; Edelman *et al.* 2004).

Using the GONG+ Doppler (line-of-sight) velocity and magnetic data-cubes taken at a cadence of 1 minute, we have evaluated the net magnetic flux and Doppler velocity during 2003, October 27/12:26–October 28/16:09 UT for the four subareas marked by the boxes in Fig. 4. Time profiles for these tracked regions are given in Fig. 4. The evolutionary trend of flux variation for the boxes ‘1’, ‘2’ and ‘3’ is found to go through a change before and after the time of the flare maximum marked by an arrow. Of course, the spikes at the flare maximum are probably caused by the changes of spectral line at that time. On the other hand, net flux for the box ‘4’, located in a flare-quiet area, shows only a gradual decreasing trend of evolution. The profiles of net Doppler velocity have larger scatter, and any flare-related changes are not obvious.

These observed flux changes may occur due to a variety of processes, such as, emergence or submergence of flux and horizontal motions leading to the enhancement or cancellation of fluxes. GONG VMG movie shows such features at several locations during the course of the flare. It is known that photospheric horizontal motions of magnetic fluxes could lead to the accumulation of magnetic energy in a flare current

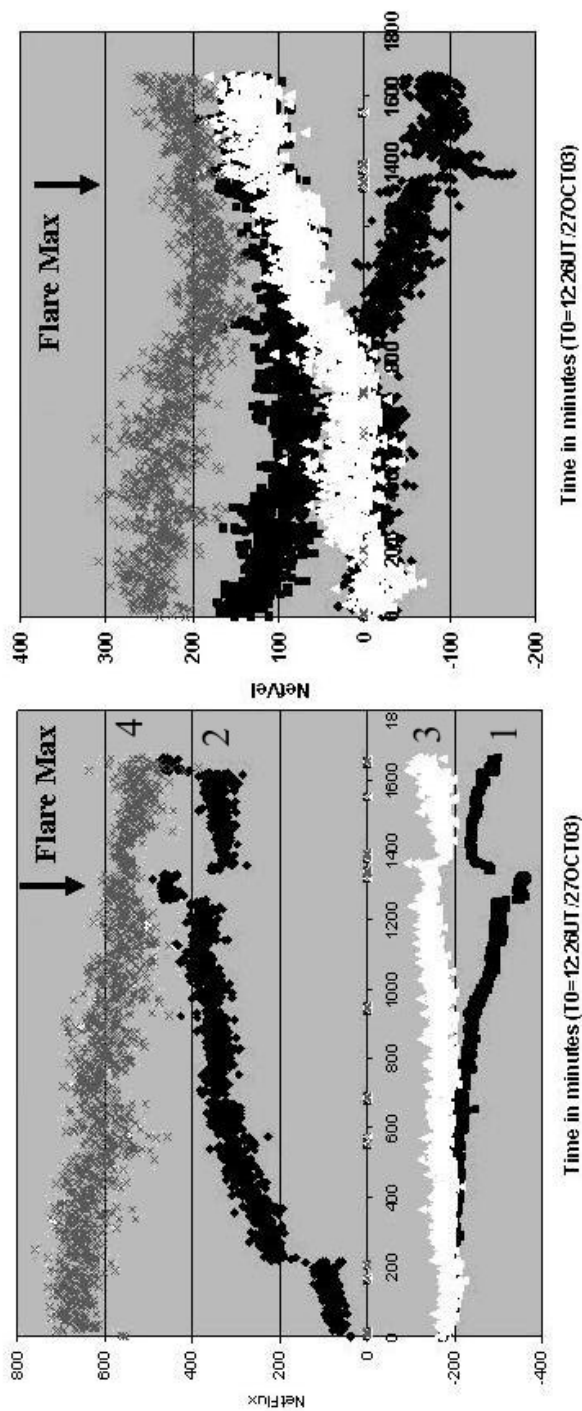


Figure 5. Temporal variations in net magnetic fluxes and Doppler velocity in the selected areas of interest.

sheet of the order of $10^{32} \sim 10^{33}$ ergs (Bilenko *et al.* 2002), and twisting of magnetic loops may result into kink instability and eventual release of 35–50% of the free magnetic energy (Gerrard *et al.* 2002). Also, magnetic modelling of flares has revealed onset of flare due to reconnection of emerging flux in a sheared magnetic field (Berlicki *et al.* 2004; Brooks *et al.* 2003). Appearance of certain horizontal velocities may lead to changes in local magnetic field structures, and generate electric current systems (Martres *et al.* 1982; Fontenla *et al.* 1995).

5. Helioseismic signatures of the flare

It is generally accepted that acoustic modes are always present on the Sun, excited by turbulence in the convection zone (Goldreich & Kumar 1988). Transient phenomena, such as flares and CMEs may contribute additional energy to these modes. However, using GONG+ time series of low-l p-mode power for the period of May 01, 1995–December 31, 2001, a mild anti-correlation is found on a global scale with disk-integrated flare-index and CME-counts (Ambastha & Antia 2004). It is interesting to ask if large flares or CMEs may have any effect on p-modes on the smaller spatial scales of active regions. Although most of the flare energy is released in the chromosphere and corona, the white light flares imply that a part of energy may be deposited in the photosphere, affecting the modes of solar oscillation as reported by Kosovichev and Zharkova (1998, 1999) and Donea *et al.* (1999). The effect of flares on solar oscillation modes has been further explored for several active regions using the ring diagram technique (c.f., Hill 1988) by obtaining 3D spectra before, during and after the flares, giving evidence of variation in mode characteristics such as frequency, width, power and asymmetry (Ambastha *et al.* 2003). At the time of 4B/X17.2 superflare, NOAA 10486 was located ideally at the disk-center and the flare was energetically substantial to look for any flare related effects on the p-mode characteristics. Figure 6 shows that frequencies and widths of the modes are not affected significantly during the three-day period around the flare, i.e., October 27–29, 2003. However, the mode-amplitude increased substantially after the large flare of October 28, 2003 (left panel). It increased further on October 29, 2003 (right panel), as a result of more flares after the superflare of October 28, 2003 (Ambastha *et al.* 2004). Recent studies have shown that subsurface flows may also lead to p-mode characteristic variations (Zhao & Kosovichev 2004; Haber *et al.* 2004).

We have used the regions tracked for 8192 minutes around the central meridian passage, and inverted the frequency differences to calculate sound speed between an active and quiet region for NOAA 10484, 10486, and 10488. This has been carried out by using the two common methods of solving the inversion integral, viz., the Regularized Least Squares (RLS) (Antia & Basu 1994), and Optimally Localized Averages (OLA) (Pijpers & Thompson 1992). These two inversion methods are complimentary in nature, therefore, one can be more confident if OLA and RLS results are in good agreement (Basu *et al.* 2004). It is found that the sound speed in active regions was lower just below the surface but at depths exceeding 7 Mm, the trend reversed. Similarly, the fitted velocities for each mode is inverted to calculate the meridional (u) and zonal (v) components as a function of depth (Ambastha *et al.* 2004). A steep gradient in meridional velocity is found to exist below a depth of 4 Mm which appears to be correlated with flares (Fig. 7) as similar a feature was also observed for other major active regions such as NOAA 9026 and 9393 (Ambastha *et al.* 2003).

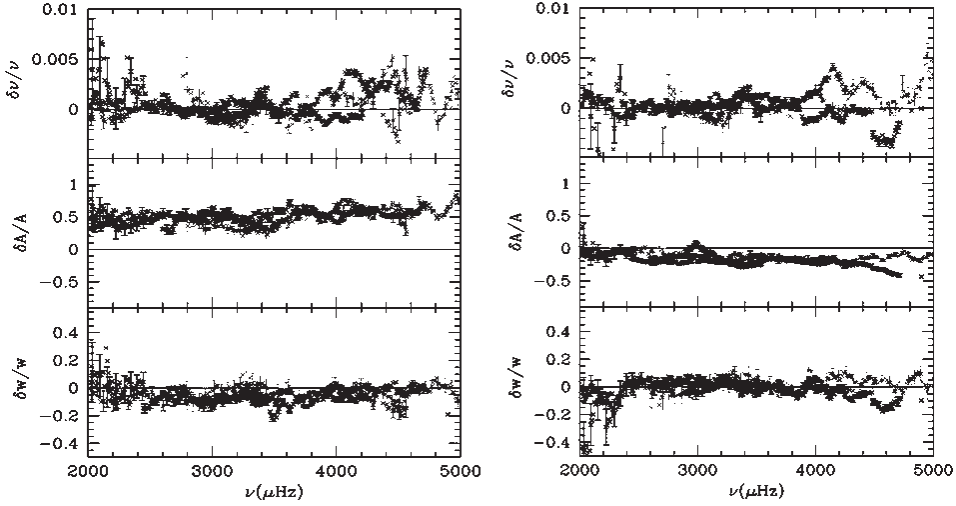


Figure 6. The relative difference in mode frequencies, widths and ratio of peak power for NOAA 10486 using ring-diagram analysis. The left and right panels respectively show the results for p-modes with $n = 0, 1, 2, 3$ obtained for the flare/pre-flare (i.e., Oct. 28/Oct. 27) and the flare/post-flare (i.e., Oct. 28/Oct. 29) phases.

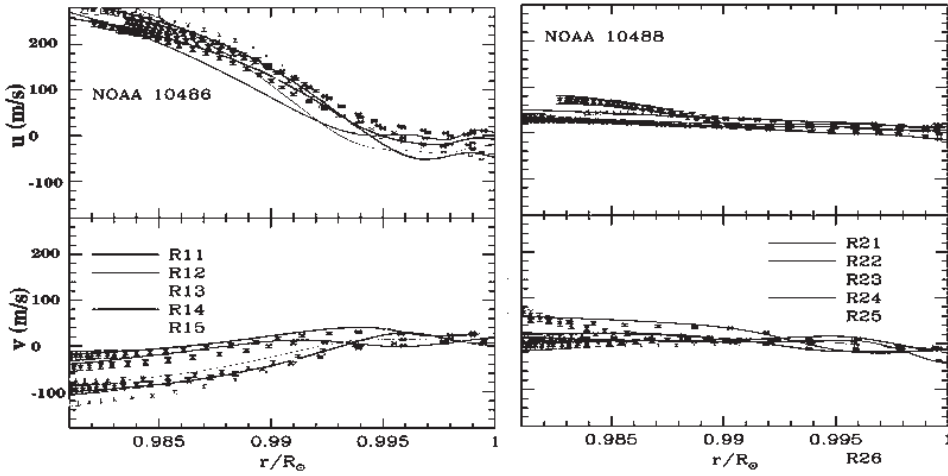


Figure 7. The meridional (u) and zonal (v) components of velocity for NOAA 10486 and 10488 at different stages. The lines show the results obtained using RLS inversions, while points mark those obtained using OLA (*c.f.*, Ambastha *et al.* 2004).

6. Conclusions

Net flux and Doppler velocity have been obtained at several locations in NOAA 10486 using GONG data during October 28–29, 2003, which show variations before and after the X17.2 flare of October 28, 2003. The Doppler maps show that upward (and downward) flows developed near the large H_α filament, associated with new

emerging fluxes. This process could have destabilized the filament delineating the narrow channel. The filament erupted along with the flare, which perhaps resulted in the fast Earth-ward moving CME. After the flare, the filament was partially restored with significant restructuring and disconnection. Evidence of reconnection was observed at the site of flux emergence prior to the flare onset. We found significant increase in the power of the acoustic modes during the X17 flare, which was beyond the normal value expected from the influence of magnetic fields. In addition, the meridional velocity in the flaring region was found to have a steep gradient below a depth of 5 Mm.

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Extremely Energetic 4B/X17.2 Flare and Associated Phenomena

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Abstract. We observed 4B/X17.2 flare in H α from super-active region NOAA 10486 at ARIES, Nainital. This is one of the largest flares of current solar cycle 23, which occurred near the Sun's center and produced extremely energetic emission almost at all wavelengths from γ -ray to radio-waves. The flare is associated with a bright/fast full-halo earth directed CME, strong type II, type III and type IV radio bursts, an intense proton event and GLE. This flare is well observed by SOHO, RHESSI and TRACE. Our H α observations show the stretching/de-twisting and eruption of helically twisted S shaped (sigmoid) filament in the south–west direction of the active region with bright shock front followed by rapid increase in intensity and area of the gigantic flare. The flare shows almost similar evolution in H α , EUV and UV. We measure the speed of H α ribbon separation and the mean value is $\sim 70 \text{ km s}^{-1}$. This is used together with photospheric magnetic field to infer a magnetic reconnection rate at three HXR sources at the flare maximum. In this paper, we also discuss the energetics of active region filament, flare and associated CME.

Key words. Sun: Flares, coronal mass ejections, reconnection.

1. Introduction

The period of late October to early November 2003 was characterized as the highest levels of solar activity to date when the solar cycle 23 was well into the decline phase. The extreme level activity began with the appearance of three most complex superactive regions NOAA 10484, 10486 and 10488. These active regions produced a series of powerful flares, large coronal mass ejections (CMEs) as well as strong space weather disturbance. NOAA 10486 grew to be the largest sunspot group of the current solar cycle and it produced many spectacular solar events, including the most energetic flares 4B/X17.2 on 28 October, the 2B/X10 on 29 October and the 3B/X28 on 4 November 2003 and their associated fast halo CMEs, different types of extraordinary radio bursts. Many authors studied these flares and associated phenomena at different wavelengths (Luthi *et al.* 2004; Pick *et al.* 2005; Klassen *et al.* 2005; Donea and Lindsey 2005; Gopalswamy *et al.* 2005; Grechnev *et al.* 2005).

In this paper, we present the preliminary analysis of X17.2 flare, observed on 28 October 2003. We examined the H α morphology of the flare and compared it with TRACE EUV/UV, RHESSI HXR and MDI magnetogram data. We also studied the energetics of sigmoid filaments, flare and associated CME.

2. Observations

During the decay phase of solar cycle 23, on 28 October 2003 we observed a long duration super flare 4B/X17.2 in $H\alpha$ from NOAA 10486 at the location S16 E08, at ARIES, Manora Peak, Nainital, India. The observations were using 15 cm, f/15 Coudé Solar Tower Telescope equipped with Bernhard Halle $H\alpha$ filter and Wright Instrument CCD camera system (16 bit, 385×576 pixel, pixel size = 22 micron square). With the help of barlow lens the image has been magnified twice, so we get a resolution of $1''$ per pixel. The observed filtergrams were corrected using dark current and flat field images taken during the observations. All images were re-registered.

The flare has been simultaneously observed by RHESSI (Lin *et al.* 2002), TRACE (Handy *et al.* 1999), SOHO/LASCO (Brueckner *et al.* 1995) and SOHO/EIT (Delaboudiniere *et al.* 1995). This flare also produced several extremely strong radio bursts and it was well observed by the IZMIRAN, Potsdam Observatory and WAVES/WIND (Bougeret *et al.* 1995) spectrometers in a broad range of radio frequencies. To understand the magnetic topology of the active region we used SOHO/MDI (Scherrer *et al.* 1995) data.

3. Analysis and results

3.1 $H\alpha$ observations

The evolution of the flare in $H\alpha$ is shown in Fig. 1. The $H\alpha$ movie shows large scale mass motions, darkening of filaments and brightening at many places in the active region before the flare onset. The first image taken well before flare at 01:28:36 UT shows the three highly twisted long sigmoid filaments (F1, F2 and F3), which started eruption before the onset of 4B/X17.2 class flare. The eruption of all three filaments continued during the flare.

The eruption of filament F1 plays an important role in the process of flare triggering and initiation of the CME and associated radio bursts. The filament F1 eruption shows detwisting in the clockwise direction. Before the major X17.2 flare an M8.4 class long duration flare peaked at 10:48 UT in the eastern side of the active region. This flare may be treated as the precursor phase for the main flare.

The 4B/X17.2 flare centered in the western part of the active region is the second largest flare event recorded during the solar cycle 23. The flare shows multi-ribbon structure in $H\alpha$. It consists of two main parallel ribbons along the sigmoid filament in opposite magnetic polarity regions (c.f. Fig. 2). These parallel $H\alpha$ ribbons show the separation as shown by a typical two ribbon flare. We have estimated the ribbon separation speed and found it very high $\sim 70 \text{ km s}^{-1}$. Between 10:25 and 10:30 UT a remote ribbon appeared in south-west direction (filaments F1 and F3 eruption direction). The chronology of the $H\alpha$ flare is given in Table 1.

To study the temporal evolution of the flare in $H\alpha$, we selected the three $H\alpha$ kernels shown in Fig. 2. The temporal evolution of these flare kernels are shown in Fig. 3. Temporal evolution of the kernels show that all kernels evolve differently. Kernel 2 has more peak period in comparison to kernel K1 and K3. The time profiles show multiple peaks during its impulsive and maximum phase. This implies that multiple magnetic field reconnection occurs in the flaring processes successively. In millimeter

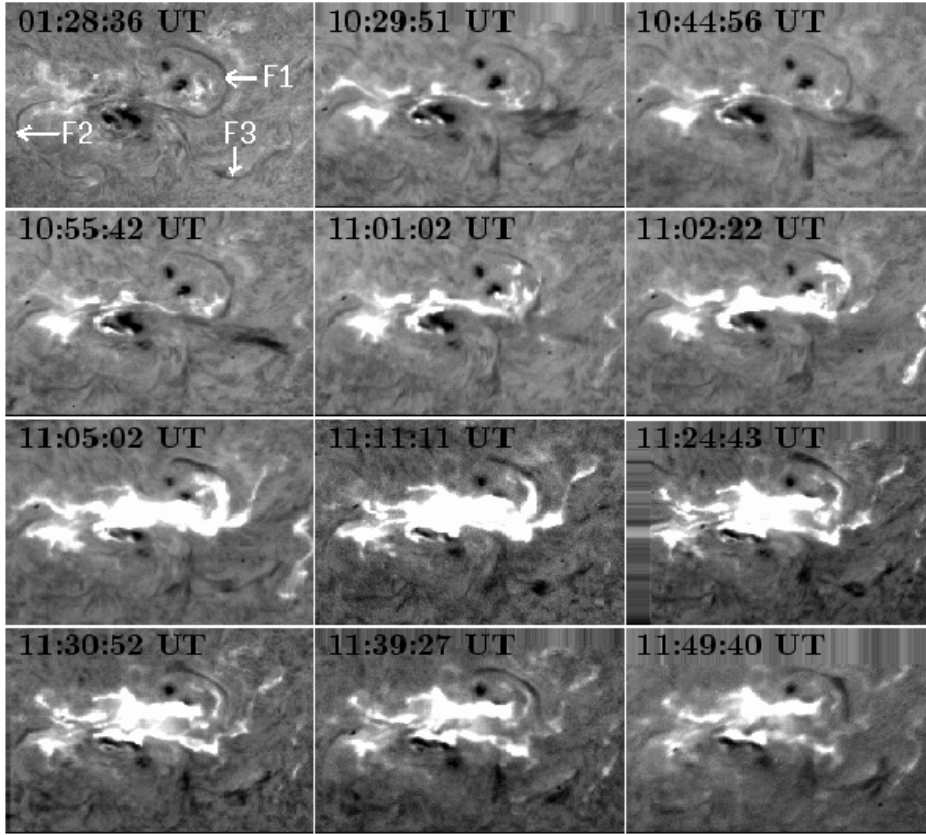


Figure 1. Evolution of historical 4B/X17.2 flare in $H\alpha$. North is up and west is to the right and the FOV is $505'' \times 384''$.

Table 1. $H\alpha$ chronology of the flare.

Time (UT)	Important changes
06:00–07:00	Activation of filaments F1, F2 and F3 started
08:22	Bright knot like structure appeared near the main flare site
09:56	M1.3 flare started in the central part of the active region
10:09	M2.3 flare started in the eastern part of the active region
10:00–10:30	Sigmoid filament eruption started in S–W direction
10:25–10:30	Remote ribbon appeared in SW direction
10:48	M8.4 long duration flare peaked
10:58	Onset of massive X17.2 flare
11:02	Remote ribbon becomes very bright
11:00–11:11	Fast development of flare ribbons, fast ribbon separation starts and reached flare attains maximum
11:11	Flare maximum phase and the remote ribbon faded out
After 11:12	Bright $H\alpha$ LOOP appeared joining kernel K1 and K3

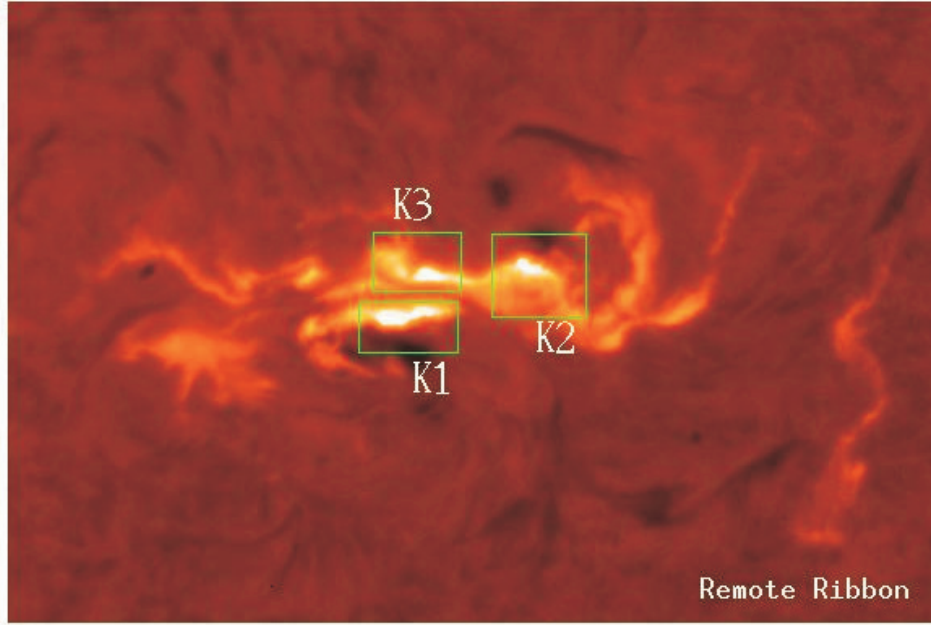


Figure 2. $H\alpha$ image at flare maximum (11:04:36 UT). The different $H\alpha$ kernels and remote brightening are shown in this figure. The FOV of the image is $574'' \times 384''$.

wavelength emission at 210 GHz shows two main features: a slowly varying component (11:02:30 to 11:2 UT) and a short lived component exhibiting three distinctive peaks between 11:03:45 and 11:06:10 UT (Luthi *et al.* 2004). A similar type of temporal evolution is also visible in $H\alpha$ kernel K2, which indicates the acceleration of high energy electrons ≥ 0.5 MeV.

3.2 Comparison of $H\alpha$ data with UV, EUV and RHESSI data

Figure 4 shows the start, maximum and decay phase of the 4B/X17.2 flare in $H\alpha$, UV (1600 Å) and EUV (195 Å) respectively. On comparison of these images we can say that the flare shows a similar type of morphological evolution in $H\alpha$, UV and EUV wavelengths. The flare ribbon's twisted structure which gets detwisted during flare evolution, is clearly visible in $H\alpha$ and UV. Evolution of post flare loops joining the parallel ribbons show that the reconnection of higher and higher loops indicate prolonged energy release and relaxation of highly sheared magnetic field.

Figure 5 illustrates the plot of the magnetic field on $H\alpha$ (upper left), MDI magnetogram overlaid by RHESSI HXR 100–200 keV sources location (marked by circle, upper right), TRACE 195 Å image overlaid by RHESSI 100–200 keV HXR contours (lower left) and the location of HXR sources at the selected times (lower right).

On comparison of the spatial location of $H\alpha$ kernels and the location of HXR sources, it appears that the $H\alpha$ kernels K1, K2 and K3 are associated with three HXR sources (c.f. Figs. 2 and 5). According to the study of Krucker & Hudson

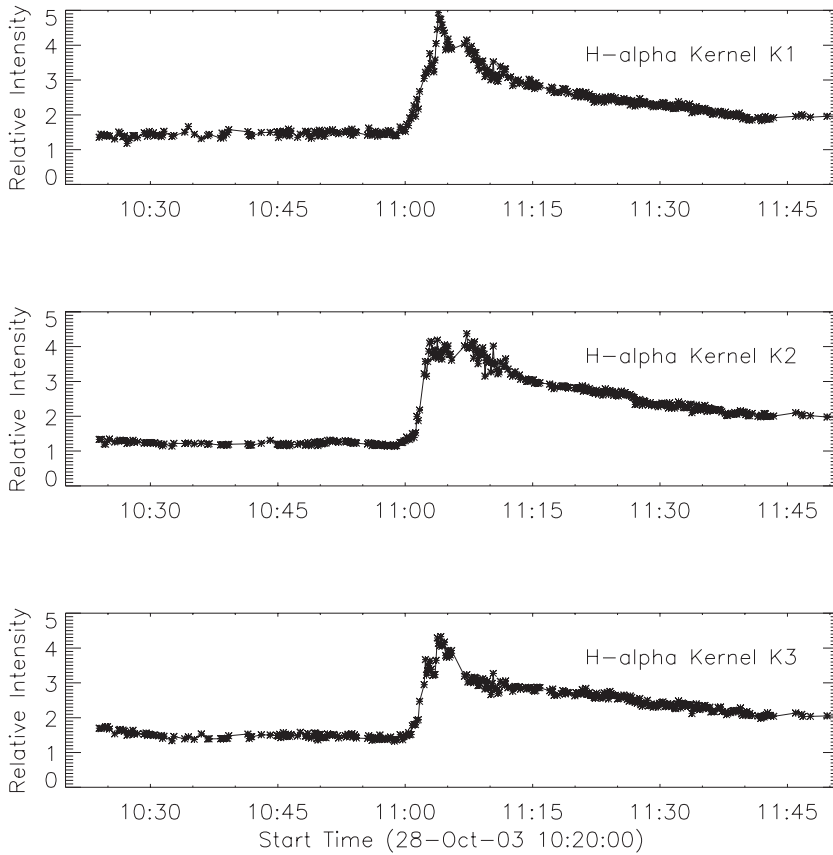


Figure 3. Temporal evolution of different flare kernels in $H\alpha$.

(2004) the two γ ray footpoints are embedded in the flare ribbons. Although close to the HXR centroid position, each γ ray footpoint is displaced to the east direction. Comparing the γ ray sources and $H\alpha$ kernel locations it seems that the γ ray sources are close to the $H\alpha$ kernels K1 and K2. Donea & Lindsey (2005) detected seismic wave emitted from this flare. The major acoustic sources are seen to coincide closely with HXR emissions from magnetic footpoints that terminate in the umbral or penumbral chromosphere of sunspots. We measure the strength of the magnetic field using MDI magnetogram at the $H\alpha$ kernels K1, K2 and K3 associated with HXR sources and the measured values are 1000, -800 and -500 gauss respectively.

Figure 5 (lower right panel) shows the position of HXR sources at some selected times (taken from Sam Krucker web page). In general the motion of HXR sources during flare depends mainly on the magnetic field configuration. During a flare, magnetic reconnection provides powerful fluxes of energy along the reconnected field lines. As the flare progresses, the newly reconnected footpoints of the magnetic field lines move away from neutral line with a velocity proportional to reconnection rate. Similar to this explanation we noticed the motion in HXR sources away from the magnetic neutral line (c.f. Fig. 5).

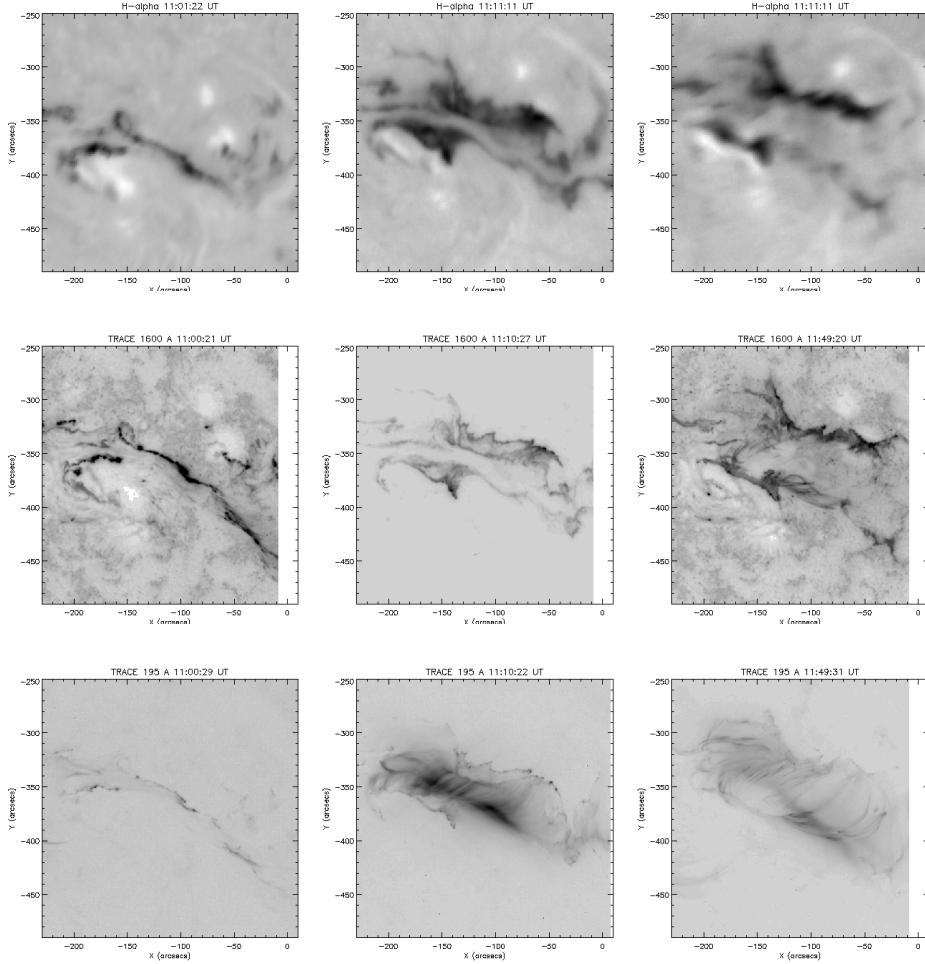


Figure 4. $H\alpha$, TRACE 1600 Å and 195 Å negative images of the 4B/X17.2 flare at begin, max and in decay phase. North is up and west is to the right.

3.3 Active region complexity

On 28 October 2003, NOAA AR 10486 was located at S16 E08 and passed over the solar disk from 23 October to 4 November 2003. The active region has $\beta\gamma\delta$ magnetic configuration. It was the largest and most complex active region of current solar cycle 23. During its passage on solar disk it produces 7 X class, 15 M class and many small solar flares. In the active region a thin corridor of negative polarity flux bisects the strong positive polarity area in the central part of the active region. The presence of sigmoid filaments in the active region shows a highly sheared magnetic field in which a large amount of free magnetic energy gets stored (Singh & Gupta 1995; Uddin & Verma 1998; Schmieder *et al.* 1993). Zhang *et al.* (2003) also reported the high shear $\sim 90^\circ$ in this active region. From the SOHO/MDI magnetogram movie we noticed the fast emergence and cancellation of magnetic flux in the active region.

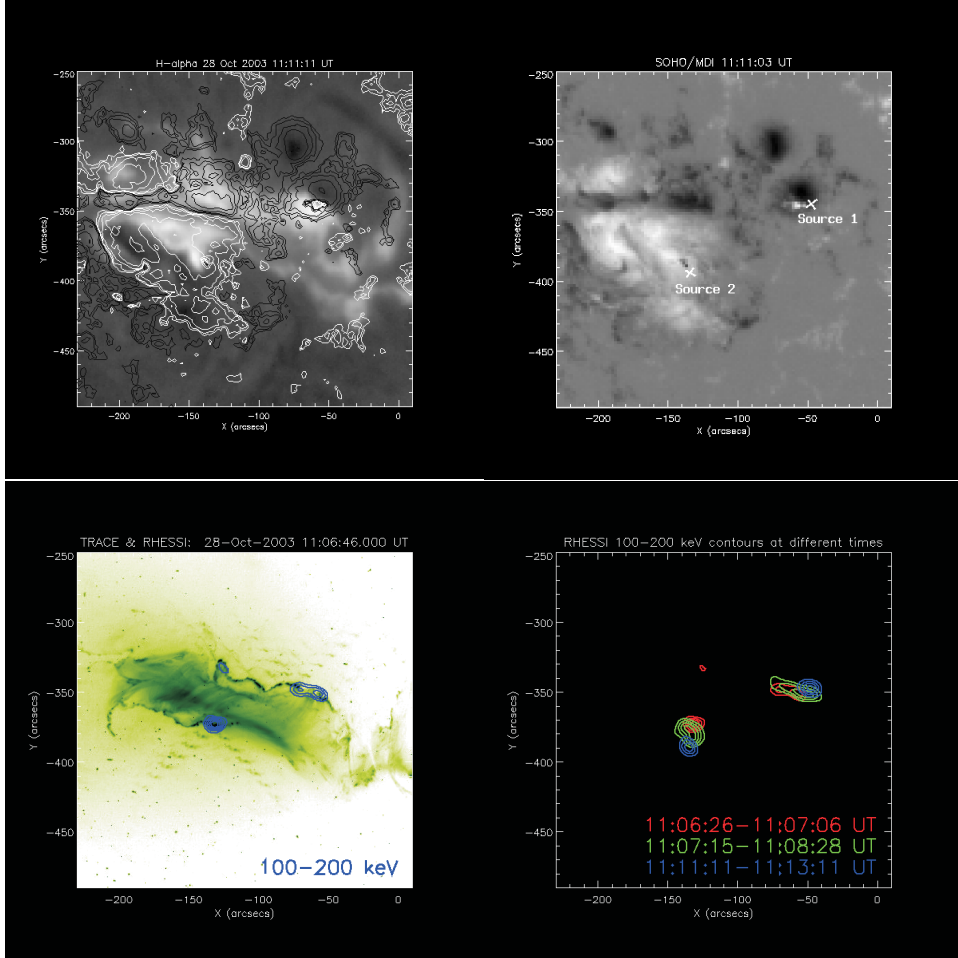


Figure 5. H α image of the flare overlaid by MDI contours (level = $\pm 100, \pm 200, \pm 400, \pm 800, \pm 1600$ gauss), MDI magnetogram overlaid by location of RHESSI HXR source (circle symbol), TRACE 195 Å image overlaid by RHESSI source and location of RHESSI HXR sources at three selected times. North is up and west is to the right.

3.4 Energetics of the filament, flare and CME

The disappearance of the sigmoid filaments is an indicator of a large scale change of coronal magnetic field (Tang *et al.* 1999). Dynamic filament eruption are generally associated with energetic solar flares, CMEs, etc. Therefore the energy stored in filament plays an important role in the production of these energetic phenomena.

Taking this into account we tried to estimate the energy stored in the filaments F1, F2 and F3, which erupted producing 4B/X17.2 superflare and high speed halo CME. The eruption of three major filaments was associated with this superflare and associated phenomena. So to understand the energetics of these filaments and its association with flare and CME energetics we estimated the total energy stored in all three filaments using the circuit model of Martens *et al.* (1986) and Martens & Quin (1989). The

parameter for the energy calculation is derived using our H α data. The estimated values of the energy stored in these three active region sigmoid filaments are 5.26×10^{34} , 1.20×10^{34} and 1.71×10^{34} ergs respectively and the total energy stored in all three filaments is 8.16×10^{34} ergs.

Following White *et al.* (2005) we measure the temperature and emission measure during flare maximum phase using the GOES data and the values are 22.24 MK and $42.14 \times 10^{49} \text{ cm}^{-3}$ respectively. Using these temperature and emission measure values we calculate the thermal energy (E_{th}) content of the flare applying the following formula:

$$E_{\text{th}} = 3 kT \sqrt{EM \cdot V},$$

where k , T , EM and V are the Boltzmann constant, temperature, emission measure and volume of the source respectively. The volume of the source is calculated using the RHESSI 10–20 keV images taking the 45% of contour level. The calculated value of the thermal energy content in the flare is 1.9×10^{31} ergs.

Magnetic reconnection is the main possible cause for the energy release process in solar flares. The reconnection rate is one of the most important parameters in this process. The reconnection rate can be calculated using the formula $\dot{\Phi} = B_c v_i$ (Asai *et al.* 2002), where B_c and v_i are the coronal magnetic field and inflow velocity into the reconnection region respectively. Assuming the magnetic field is conserved then the reconnection rate can be expressed as $\dot{\Phi} = B_p v_f$, where B_p is the photospheric magnetic field and v_f is the ribbon separation speed respectively. For this flare we calculated the value of the reconnection rate at three HXR sources which are spatially correlated with H α kernels viz., K1, K2 and K3. The calculated values of the reconnection rate at kernels K1, K2 and K are 70 volt cm^{-1} , 56 volt cm^{-1} and 35 volt cm^{-1} respectively.

The mass, projected velocity and kinetic energy of the CME associated with this flare have been given in SOHO LASCO catalog (http://cdaw.gsfc.nasa.gov/CME_list/UNIVERSAL/text_ver/). The values of the mass, projected velocity and the kinetic energy of the CME are $4.0 \times 10^{16} \text{ gm}$, 2459 km s^{-1} , $1.2 \times 10^{33} \text{ erg}$ respectively. We compared the value of kinetic energy and the value of the thermal energy content of the flare and found that the thermal energy content of the flare is 2 orders smaller than the kinetic energy of the CME.

3.5 Coronal mass ejections and other associated phenomena

A full halo CME with very bright emittance all around the occulting disc of the SOHO/LASCO coronagraph was associated with this flare. The event was first observed on C2 at 10:54 UT as a bright loop front over the West limb, by 11:30 UT the front had developed into a full halo CME. The front first appeared as a full halo CME in C3 at 11:42 UT. In the EUV region, according to SOHO/EIT data, it was accompanied by a coronal wave propagating in the northern hemisphere and dimmings extended over all the southern hemisphere. The CME was very fast and the speed was 2458 km s^{-1} . Comparing the direction of filaments (F1 and F3) eruption and the CME direction indicates that the CME was associated with the eruption of these filaments.

The flare was associated with several different types of radio bursts. The radio bursts consist of a group of rapidly drifting type III bursts and a complicated system of more

slowly drifting diffuse type II (speed 1250 km s^{-1}) with fundamental-harmonic structure followed by a type IV continuum. This flare has intense radio bursts including a 245 MHz burst near 500,000 sfu, a Tenflare of 13,000 sfu. The space weather disturbances associated with this flare are characterized by an extremely strong increase in the high-energy proton flux ($29,500 \text{ pfu} > 10 \text{ MeV}$) and a severe geomagnetic storm with D_{st} index -363 nT . This flare is a strong cosmic ray event and produces unusual Ground Level Enhancement (GLE). Bieber *et al.* (2005) analyzed and described many unusual features of the GLE associated with this flare.

4. Discussion and conclusions

The 4B/X17.2 flare observed at ARIES in $H\alpha$ was associated with filament eruptions, fast full halo CME, different types of strong radio bursts, strong proton flare, etc.

The flare shows the multi-ribbon structure. The flare has two main parallel ribbons in the opposite magnetic polarity regions. From the comparison of $H\alpha$ flare in UV and EUV wavebands we conclude that the flare shows a similar type of morphology in $H\alpha$, EUV and UV wavelengths.

We measure the value of the magnetic field at three HXR sources and find that it corresponds to high magnetic field strength. Normally the HXR sources are found to be associated with weak magnetic field regions. However, Asai *et al.* (2002) found the HXR sources to be associated with high magnetic field regions. So our results confirm the finding of Asai *et al.* (2002).

We derived the magnetic reconnection rate at the HXR sources measuring the $H\alpha$ ribbon motions and photospheric magnetic field at HXR sources. We found that reconnection rate at kernels K1, K2 and K3 are 70 volt cm^{-1} , 56 volt cm^{-1} and 35 volt cm^{-1} respectively. The values are one order greater than the reported value of Poletto and Kopp (1986) and comparable with the values of Qiu *et al.* (2002). Such large values of magnetic reconnection rate may be due to the fact that the flare kernels are moving towards the sunspots and cover the sunspot umbrae, where the magnetic field is very strong.

We compared the kinetic energy of the CME, thermal energy content of the flare and stored energy in the twisted sigmoid filaments (F1, F2 and F3). The thermal energy content of the flare was found to be 2nd order of magnitude less than the CME kinetic energy. Recently Emsile *et al.* (2004) analyzed two events, and found that the CME kinetic energy was found to be almost 1 order larger than the flare energy. We also compared the energy stored in sigmoid filaments, flare thermal energy and the kinetic energy of the CME and found that the energy stored in the sigmoid filament is 1 order higher than CME energy, however 3 orders higher than the thermal energy content of the flare. Hence we conclude that the filaments F1, F2 and F3 have more than sufficient energy to trigger the flare, CME and other energetic phenomena.

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The Flares Associated with the Dynamics of the Sunspots

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Abstract. In the present study, we consider six years data of spot groups that have well developed leading and following spots obtained from the Kodaikanal Observatory white light pictures and occurrence of H α flares. From the daily observations, we compute the variations in rotation rates, meridional velocity, the areas and longitudinal separations. We find that among all these variations, the occurrence of abnormal rotation rates (the rotation rates that have greater than 1σ) and longitudinal minimum separation during the course of their evolution eventually lead to triggering of flares. We also find that the events of abnormal rotation rates, longitudinal minimum separation and the flares occur mainly during the 50–80% of the sunspots' life span indicating magnetic reconnection probably below ($0.935 R_{\odot}$) the solar surface. Relevance of these results with the conventional theory of magnetic reconnection is briefly discussed.

Key words. Sunspots — rotation — magnetic flux — magnetic reconnection.

1. Introduction

Sunspots are supposed to be associated with many transient activity phenomena such as flares, coronal mass ejections (CME), etc. The following studies show the association between changes in the sunspots' dynamics, emerging flux region, twisting of the field lines above the solar surface, magnetic shear and all of these activities together with eventual triggering of the flares.

Tanaka and Nakagawa (1973) have shown that the required energy storage for the eventual triggering of the flares is due to proper motion of the sunspots. Using disc passage observations of sunspot activity NOAA 2372 during April 4–13, 1980 Ambastha and Bhatnagar (1988) show that all the sunspots within an activity group experienced the large proper motions before the major flare events. From the analysis of polarity distributions in sunspot groups that produced solar proton flares, Sakurai (1976) inferred that the origin of solar flares may be closely related with the instability that is associated with twisting of sunspot field lines in the chromosphere and the lower corona. Ishii *et al.* (1998, 2003) using a variety of solar activity phenomena conclude that emergence of twisted flux bundle is the key cause for production of the flares. Using six year (1969–1974) observations of Kodaikanal Observatory white light pictures, Hiremath & Suryanarayana (2003) showed that either leading or following or both of the bipolar spot groups that have abnormal rotation rates during the course

of their evolution are strongly associated with the occurrence of flares in the later stage of their life span.

Another important precursor is the magnetic shear (Hagyard *et al.* 1984) that triggers the flares. The continued evolution of the magnetic flux causes the field's maximum shear to exceed a critical value of shear resulting in triggering of the flares. Later studies (Venkatakrishnan *et al.* 1989; Ambastha *et al.* 1993; Schmieder *et al.* 1997) also confirm the view that magnetic shear is one of the important properties for production of the flares.

Yet it is still an unsettled problem whether flares and CMEs are triggered by changes in the erupting magnetic flux resulting in shear, or dynamics of the sunspots or triggered by both (Schmieder *et al.* 1994) of these activities.

The conventional belief (Priest 1981; Haisch & Strong 1991; Parker 1994; Lin *et al.* 2003) is that the source of energy for producing the flares is due to *magnetic reconnection* of the opposite polarities wherein the flux is annihilated in a compact region. If one accepts this conventional belief, however, there are no conclusive evidences whether reconnection of the magnetic field lines occur in the solar atmosphere or below the photosphere. There are some studies (Zirin & Liggett 1987; Leka 1996; Canfield & Pevtsov 2000; Hiremath & Suryanarayana 2003) suggesting that reconnection events may be taking place below the photosphere. In both the cases of reconnection events that may be occurring above or below the photosphere, approaching of sunspots towards each other and then merging of foot points is necessary. The other two crucial conditions for magnetic reconnection are as follows: For magnetic field of strength of B and length scale L , energy released due to magnetic reconnection of opposite polarity field lines is $\sim L^3 B^2$. That means, in order to produce the observed flare energy ($\sim 10^{27}$ – 10^{33} ergs) (Metcalf, Leka and Mickey 2005) the length scale of reconnection region should be $\sim 10^5$ – 10^9 cms. The third crucial condition for the events of magnetic reconnection is that the inflow velocity in the vicinity of reconnecting region is ~ 0.1 times the Alfvén velocity (Petschek 1964), where Alfvén velocity is computed from the ambient magnetic field and density structure.

In order to confirm these three crucial conditions for the magnetic reconnection, using Kodaikanal Observatory data, we search for such events and show that whenever the sunspot groups (containing the leader and the follower) have the events of abnormal rotation rates and the minimum longitudinal separation, flares eventually trigger in that region. In section 2, we present the motivation for this study. The data analysis and the results are presented in section 3. The results with conclusive evidences for the magnetic reconnection are presented in the last section.

2. Motivation for the present study

The present general consensus is that sunspots originate below the solar surface due to an unknown dynamo mechanism. Due to very high conductivity of the solar plasma and assuming that rising flux tube does not acquire extra flux from the ambient medium, sunspots isorotate with the internal plasma and due to buoyancy rise towards the surface along the path of rotational isocontours. This implies that sunspots are very good tracers of the internal dynamics and structure of the solar interior. Hence if the sunspots which have first and second days appearance on the surface and if one computes their initial rotation rates, then one can infer rotation rate of the internal solar

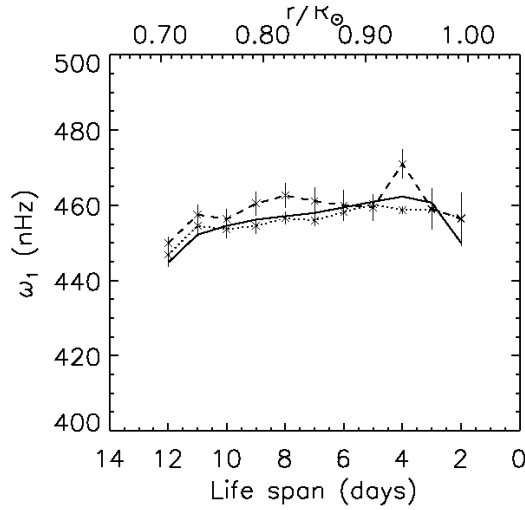


Figure 1. Initial rotation rate of the sunspot groups with respect to their life span. Continuous curve is helioseismologically inferred radial variation of rotation. Dotted curve represents the rotation rates of spot groups having areas < 100 mh and dashed curve represents the rotation rates of spot groups having areas > 100 mh. In the same plot, radius values are plotted along the top x axis.

plasma where the sunspots' foot points are anchored. Recent studies (Hiremath 2002 and references there in; Zuccarello and Zappala 2003) show that variation of the initial rotation rates obtained from the daily motion of sunspot groups with respect to their life spans is almost similar to the radial variation of the internal rotation profile of the solar plasma. From Hiremath's (2002) paper, we reproduce the results in Fig. 1 that illustrate the comparison between the variation of initial rotation rates of the sunspot groups for different life spans and radial variation of internal rotation profile as inferred (Antia, Basu and Chitre 1998) from the helioseismology. Note the striking similarity between these two profiles. In order to reach closer to the reality of the physics of convection zone and dynamics of the flux tubes, in the same study, the rate of change of initial rotation rates of the sunspot groups (that represent the acceleration or deceleration of the flux tubes in the ambient plasma) are compared (see the Fig. 5(b) of Hiremath (2002)) with the radial profile of gradient of rotation (that is computed from the radial variation of rotation of the plasma inferred from the helioseismology). Again we get a striking similarity between these profiles.

To conclude from that study (Hiremath 2002), for different life spans, initial sunspot dynamics over the surface represents the internal dynamics in different layers of the convection zone.

For example initial anchoring of a flux tube whose life span is 10 days is near base of the convection zone and initial anchoring of a flux tube whose life span is 5 days is in the middle of the convective envelope. However, steady rise of the flux tubes (that have leader and the follower) along the isorotation contours does not guarantee that their foot point separation remains constant. Owing to strong rotational shear (around $0.935 R_{\odot}$) (see the helioseismic inferred rotational results of Antia, Basu and Chitre 1998) near the surface, either or both of the opposite polarity foot points of the flux

tubes are accelerated or decelerated towards each other. That is if one computes the rotation rates, either leader or follower or both may have *abnormal rotation rates* (Hiremath and Suryanarayana 2003) during a particular life span of the sunspot group resulting in merging and reconnection (of part) of the magnetic field lines of the foot points that eventually trigger the flares. The foot points can also be merged due to meridional circulation and flux emergence in the interior that results in changing of the area of the sunspot group. Using Kodaikanal Observatory data of the sunspot groups (that have the leader and the follower spots) we test these concepts in the following analysis.

3. Data analysis and results

For the years 1969–74, we use both the data set of positional measurements (heliographic latitude and longitude from the central meridian) of the sunspot groups (that have leading and following sunspots) taken from daily white light images and the flare events in the $H\alpha$ images from the Kodaikanal Observatory. The details of the telescope and observations of daily white light images are given by Sivaraman *et al.* (1992). Using similar criteria (Hiremath 2002) in selecting the sunspot groups, we compute rotation rate ω_i of the leading and following sunspots as follows:

$$\omega_i = \frac{(l_{i+1} - l_i)}{(t_{i+1} - t_i)}, \quad (1)$$

where l is the heliographic longitude from the central meridian, t is the time of observation, $i = 1, 2, 3, \dots, n$, and n is the age of the spot group.

We compute daily *longitudinal separations* $d_i = l_L - l_F$ (l_L and l_F are the longitudes of the *leader* and the *follower*) of the foot points of the spots. Following equation (1), we compute the *rate of change of longitudinal separation* S_i

$$S_i = \frac{(d_{i+1} - d_i)}{(t_{i+1} - t_i)}. \quad (2)$$

In the following analysis we use the combined data (1969–74) for the whole region of heliographic latitudes of 0° to 40° in both the solar hemispheres. We define *life span* of a spot group as the total number of days between the first and the last appearance on the same part of the solar disk satisfying the aforementioned criteria.

For the period of observations, we select 57 well-developed spot groups that have leader and follower spots. Using equation (1) and (2), we compute daily rotation rates ω_i and rate of change of longitudinal separation S_i . Similarly we compute the meridional velocity and rate of change of areas of the sunspot groups. We find that the last two parameters are not correlating with the events of triggering of the flares and are not presented.

In Fig. 2, we present the normalized values of daily rotation rates and the rate of change of longitudinal separation observed on 28 January 1970. The continuous vertical lines are occurrence dates of the flares. The occurrence dates of flares are considered during the maximum phase. The normalized values are defined as follows: If x_i are the data points for different i days, $\bar{x}$ is the average of all data points and σ is the standard deviation of rotation rates of the leading and the following spots and the rate of change of separation of their foot points, then the normalized value is $y_i = (x_i - \bar{x})/\sigma$.

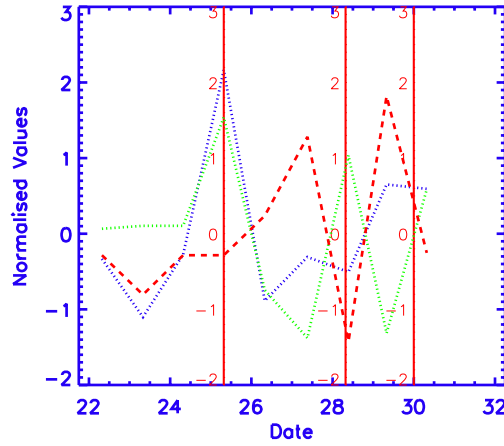


Figure 2. The normalized rotation rates and rate of change of longitudinal separation of the leader and the follower during their evolutionary phases. The blue and green dotted lines represent the normalized rotation rates of the leading and the following spots. The red dashed line represents the normalized rate of change of longitudinal separation of the leader and the follower spots. The red vertical continuous lines are the occurrence dates of the flares.

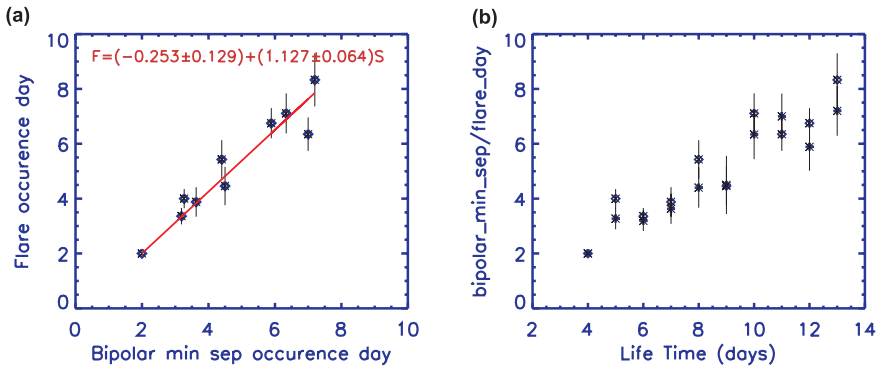


Figure 3. (a): The association between the occurrence of the minimum separation and the flare during the evolution of the spots. The red continuous line is obtained from the linear least square fit. Here S and F represent occurrence of minimum separation and the flares respectively. (b): Days of minimum separation and the flares during the evolution of spots. The symbols $\diamond$ and the square both in blue represent the day of minimum separation and the flares respectively.

Since we want to present all three parameters (rotation rates of the leader and the follower and the rate of change of longitudinal separation) that have different ranges of magnitudes, the normalization allows presentation of the three variables in a single plot. Whenever there are minimum approaching distances (represented by the negative values of the variation of separation) between the foot points of the leading and the following spots, on the same day or later the spots experience abnormal rotation rates leading to triggering of flares.

For the 57 spot groups, we note the occurrence of longitudinal minimum separation and the corresponding occurrence of the flare. The resulting correlative analysis is presented in the scatter diagram of Fig. 3(a) (left).

The correlation coefficient is found to be 94% with a very high significance ($\sim 100\%$). One has to be cautious in interpreting the magnitudes of very high correlation coefficients (~ 1). In the present analysis we compute the Spearman Rank-Order correlation coefficient and its significance (Press *et al.* 1992). This method of finding the correlation between two variabilities is more robust than the usual method (i.e., by linear correlation). From this method, we not only find a very high correlation but also at very high significance.

In order to know at what stage of a sunspot's life span the events of minimum separation and flares occur, we separate spot groups of different life spans. In Fig. 3(b) (right), we present the results with life span along the x axis and the corresponding occurrence of the minimum separation and the flares along the y axis. The errors are determined using the formula $\sigma/(N)^{1/2}$, where N is the total number of events of minimum separation and flares and σ is the standard deviation. As we found in the previous study (Hiremath and Suryanarayana 2003), for the events with abnormal rotation rates, a spot with a 4 day life span experiences on average a minimum separation and correspondingly the occurrence of a flare on the second day. A spot with a life span of six days experiences the same events on the third day and so on. In other words, abnormal rotation rates of the spots and the minimum distances of the foot points on average occur at between 50 and 80% of the life span during the course of their evolution, probably indicating annihilation of magnetic energy below the surface (Hiremath and Suryanarayana 2003).

If we assume that the flares occur due to magnetic reconnection, then it is interesting to know the magnitude of minimum separation during the occurrence of the flare. In Fig. 4(a), we present the minimum separation (in degrees) of the leading and the following foot points during the occurrence of the flare. For 57 spot groups, totally we have 85 flares and in each class we have the following number of flares: (i) sf : 1, (ii) sn : 30, (iii) sb : 8, (iv) lf : 2, (v) ln : 23, (vi) lb : 10, (vii) 2f : 0, (viii) 2n : 3, (ix) 2b : 8. Though we have less number of flares for the classes sf and lf, we have the following important conclusion. In order that reconnection events occur below the surface, the approaching spots that experience abnormal rotation

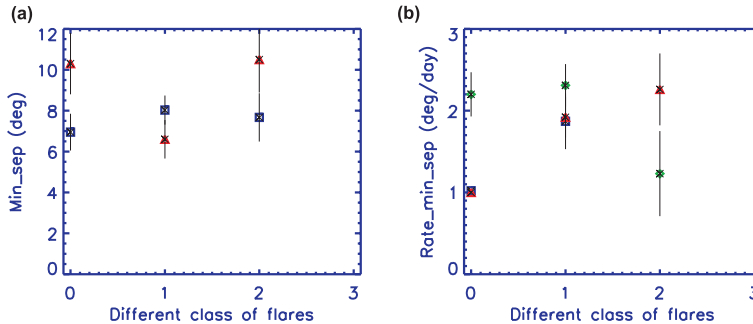


Figure 4. (a): Minimum separation for the different classes of flares: the square in blue color represents n (normal), the Δ with red color represents b (bright). Here 0 along the x axis represents the S subclass flare. The numbers 1, 2, 3 are higher subclass flares. (b): The rate of change of longitudinal minimum separation for different classes of flares: the square in blue color represents f (faint), the $\diamond$ with green color is n (normal) and the Δ with red color represents b (bright). Here 0 along the x axis represents the S subclass flare. The numbers 1, 2, 3 are higher subclass flares.

rates should have a minimum longitudinal separation, on average 6° – 10° in the photosphere. It is also interesting to know the speed at which foot points of the spots approach each other during the occurrence of the flare. In Fig. 4(b), we present the rate of change of minimum separation for different classes of flares. The foot points of the spots that eventually trigger the flares approach each other on average at a rate of $\sim 1^\circ$ – 2° /day.

4. Discussion and conclusions

Since the majority of spots that have leading and following parts are bipolar (Zirin 1988), we assume that all the spot groups that are considered in the present study are *bipolar*. Thus we can invoke the theory of *magnetic reconnection* for the interpretation of the results.

As pointed in the introduction, from the present study, we satisfy the three crucial conditions for the magnetic reconnection as follows.

As for satisfying the first condition of magnetic reconnection, Fig. 2 indicates that whenever foot points of the bipolar spots approach each other flares trigger.

To satisfy the second requirement, the present analysis (see Fig. 4(a)) shows that on the surface the average minimum separation between bipolar spots is $\sim 6^\circ$ – 10° in longitude ($\sim 10^9$ cms) during the occurrence of flare events. By taking a clue from our previous study (Hiremath and Suryanarayana 2003) that the reconnection may be occurring below the surface at a depth of $0.935 R_\odot$, from simple plane trigonometry, one can estimate the thickness (length) of the reconnecting region to be $\sim 10^5$ cms which is in the required range of 10^5 – 10^8 cms.

For satisfying the last requirement of magnetic reconnection, the strength of the background magnetic field in the vicinity of the reconnecting region is required. The region outside the sunspot has a background magnetic field strength of ~ 1 G (Stenflo 1994). This is not the same as the strength of the magnetic field (~ 40 G) of the localized small scale magnetic structures as determined by the Hanle method. On the other hand, we want to determine the strength of the large-scale global magnetic field in the sunspot-free region. Observational (Duvall *et al.* 1979; Stenflo 1994) and theoretical (Hiremath and Gokhale 1995) estimates of the magnetic field strength of such a region show that it is ~ 1 Gauss.

Thus, at the surface of the photosphere, in the region outside the sunspot, the Alfvén velocity $v_a (= B/(4\pi\rho)^{1/2})$, where B is the strength of the magnetic field and ρ is the density) is found to be $\sim 10^5$ cms/sec. The results from Fig. 4 show that the leading and the following spots that approach each other during the occurrence of the flare have a separation velocity of $\sim 1^\circ$ /day (10^4 cms/sec). This result satisfies the requirement that $v_{in} = 0.1 v_a$ (where v_{in} is the inflow velocity in the vicinity of the region of flare occurrence). *Thus, this study strengthens the conventional view that flares may be occurring due to magnetic reconnection.*

The overall conclusion of the present study is that during the course of the evolution of leading and following sunspots and in order to trigger flares, the foot points associated with the abnormal rotation rates of the leading and following spots should have an approaching velocity of 1–2 deg/day and ultimately reach a minimum separation of $\sim 6^\circ$ – 10° for probable magnetic reconnection below the surface.

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Design and Status of Solar Vector Magnetograph (SVM-I) at Udaipur Solar Observatory

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Abstract. We present the status of the instrument called Solar Vector Magnetograph Phase-I (SVM-I) currently being developed at Udaipur Solar Observatory. SVM-I is an instrument which aims to determine the magnetic field vector in the solar atmosphere by measuring Zeeman induced polarization across the spectral line. The instrument is currently in a preliminary development stage, with all components under an evaluation process. The integration of components is being done. The integrated performance of the system on a tracking mount and its control software is being tested. Some test observations of sunspots has been carried out. In this paper we give a technical description of the hardware and software elements of the instrument and present the polarized images obtained during test observations.

Key words. Sun: polarimetry—Fabry–Perot—solar magnetic fields.

1. Introduction

It is widely accepted that the source of explosive phenomena on the Sun is the interplay between magnetic fields and solar plasma. The measurement of vector magnetic fields is of great importance in order to develop a better understanding of these explosive phenomena and to be able to predict these events. The spectral lines from regions permeated by magnetic fields are Zeeman split. In visible wavelengths, the Zeeman splitting is measurable only for strong field regions like in sunspot umbrae, while, in the infra-red wavelengths one is able to observe splittings for most of the active region field strengths. However, even when the splitting is not resolved, the polarization of the Zeeman components provides a diagnostic tool for inferring the magnetic field vector. Thus, the problem of quantifying the magnetic field reduces to measurement of the full polarization state of the Zeeman components, which is represented in terms of Stokes vector.

Conventionally, the polarimetry techniques employ wave retarders for modulation like crystal waveplates, electro-optic crystals like KDP or nematic and ferro-electric liquid crystals or Photo Elastic Modulators (PEMs) with polaroids or calcites as the analyzers. For spectral isolation one uses either dispersive methods like grating spectrograph or interferometric methods like Michelson cubes or Fabry–Perot interferometers or bi-refrigrant filters.

In this paper we describe the design and status of Solar Vector Magnetograph Phase-I (SVM-I). The paper is organized as follows. The optical design and a brief layout of various components of the instrument is described in section 1. In section 2, we describe the characteristics of various subsystems of the magnetograph. Section 3 describes the control software used for acquiring data and controlling the instrument. In the last section, the first-light images of a sunspot and very preliminary analysis is presented. The evaluation of the data is under progress, which will eventually drive the design modifications in second phase of solar vector magnetograph.

2. Schematic description of the instrument

2.1 Imaging system

Figure 1 shows the schematic layout of the instrument. The primary imaging is done using a Schmidt–Cassegrain telescope of 20 cm aperture. A broadband interference filter of bandpass 15 nm is placed in front of the telescope to prevent excessive heating of the telescope optics. The bandpass is centered around 630 nm to allow the Fe I line pair in which we are interested to do polarimetry. The focal length of the telescope is 2 metres, this makes an $f/10$ beam. The image of the Sun is formed on a field stop where a circular aperture selects a desired field-of-view on the Sun.

2.2 Polarimeter

The light passing through the aperture is analyzed for polarization using the polarimeter module which consists of two quarter wave-plates mounted on a high resolution rotary stage. These wave-plates, which are compound zero-order quartz wave-plates, are used

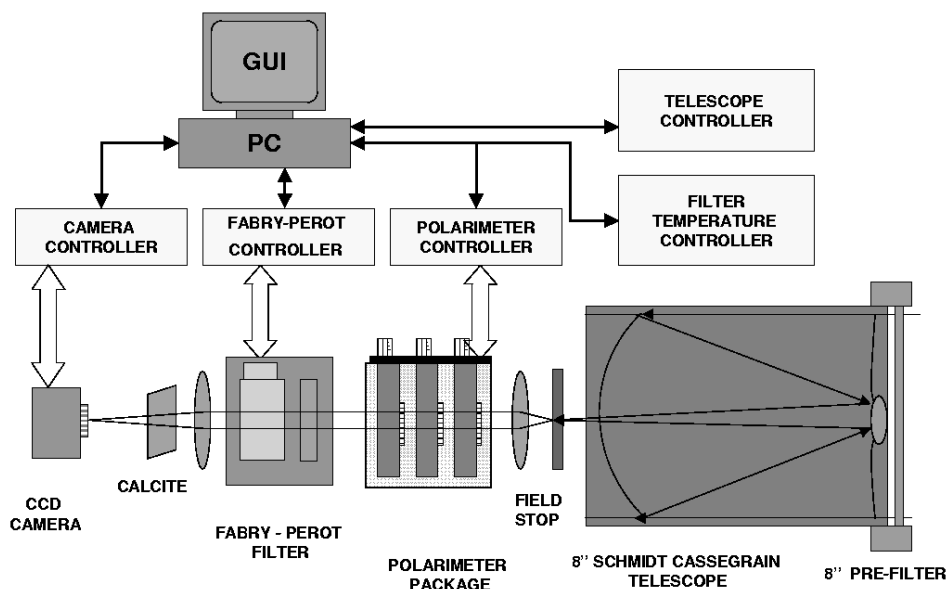


Figure 1. A schematic layout representing the optical and controls layout of the Solar Vector Magnetograph.

Table 1. Wave-plate orientations for the polarimeter.

Stokes parameter	Wave plate 1	Wave plate 2
$I \pm Q$	0°	0°
$I \mp Q$	45°	45°
$I \pm U$	0°	45°
$I \mp U$	0°	135°
$I \pm V$	45°	0°
$I \mp V$	135°	0°

to modulate the polarized light signals. The retardation properties of these wave-plates were tested in a laboratory using a technique which employs Babinet compensator and CCD camera, as described in Sankarasubramanian & Venkatakrishnan (1998). The orientation of the wave-plates are adjusted through computer control to filter out desired Stokes parameters. The orientations of the two wave-plates for different Stokes measurement are given in Table 1. A total of three measurement positions are required to measure the complete Stokes vector. Since, the polarimetry is done in diverging beam and retardation of wave-plates vary with angle-of-incidence, this has two effects. Firstly, the retardation varies across the field-of-view (FOV), which is negligible as our FOV is only few arc-minutes. Secondly, the value of retardation has a spread around a mean value. The amount of spread depends on half of the cone-angle of the diverging beam. In our case for $f/10$ beam the half cone-angle is 2.8° , which corresponds to a variation of retardation of around 1.5° from true retardation value, which is not significant considering the errors of retardation across the area of retarder and misalignment errors in polarimeters. In the present design the retarders are not kept in collimated beam because there will be a systematic variation of retardance across the field-of-view and also the reflections from the back surface of wave-plates produces problematic ghost images. The analysis of polarization is done using two crossed-calcite polarization analyzer assembly placed just before the CCD camera.

2.3 Fabry–Perot filter & CCD camera

After the polarization modulation by the wave-plates, the beam is collimated using a lens of focal length $f = 180$ mm. This amplifies the angles of the incoming light by about a factor of 11. The collimated beam passes through Fabry–Perot (FP) etalon. The etalon is an air spaced peizo-scanned and servo controlled etalon (Hicks *et al.* 1984). The FWHM of the transmission band of FP is determined to be about $120 \text{ m}\text{\AA}$ using laser in laboratory (Choudhary & Gosain 2002) and also using telluric lines formed in Earth's atmosphere. The collimated design suffers from the problem of spectral shift across the field-of-view but offers better spectral resolution than telecentric design and is more compact. The scanning FP etalon can be calibrated for FOV shift as full line profiles are available. The adjacent transmission peaks of the FP are blocked using a narrow-band interference filter. This filter is a two-cavity filter with a pass-band of 0.35 nm . The filter is kept in a temperature controlled oven with temperature stability of $\pm 0.5^\circ\text{C}$. The beam is finally re-imaged using a lens of $f = 500$ mm which gives a magnification of 2.7 times. The final image scale is around 0.87 arcsec per pixel. The converging beam is passed through a calcite polarization analyzer assembly placed

just before the CCD camera. This makes two identical images of the region in orthogonally polarized light. The two images are simultaneous and hence immune to seeing induced spurious polarization signals. The CCD camera is Apogee AP6E camera, which uses a KAF 1001E image sensor consisting of 1024×1024 pixel array with square pixel of size $24 \mu\text{m}$. The pixels are digitized up to 14 bits at a speed of 1 MHz. Typical readout time for full frame, without binning, is about 1.3 seconds.

2.4 Tracking system and housing

The entire assembly of SVM optics is straight and symmetric about the optic axis with no oblique reflections. Thus, the instrumental polarization is expected to be minimal. The stress induced polarization in the broad-band eight inch pre-filter is yet to be determined, but is expected to be negligible as was the case in the MSFC magnetograph (Hagyard *et al.* 1982). The length of the SVM optical setup is about 1.6 meters in length. The individual modules are fixed on to movable base plates, which are then fixed on optical rail at the desired locations. The entire rail is mounted on German Equatorial mount. Presently the mount works without solar guider, but the polar alignment is accurate enough to track a sunspot within a few arc-seconds for a time duration of 5 to 6 minutes. This duration is long compared to the time required to make one complete Stokes measurement. The entire system is housed on the roof-top of USO main campus, a temporary platform with a sliding dome is developed to house the instrument (see Fig. 2).

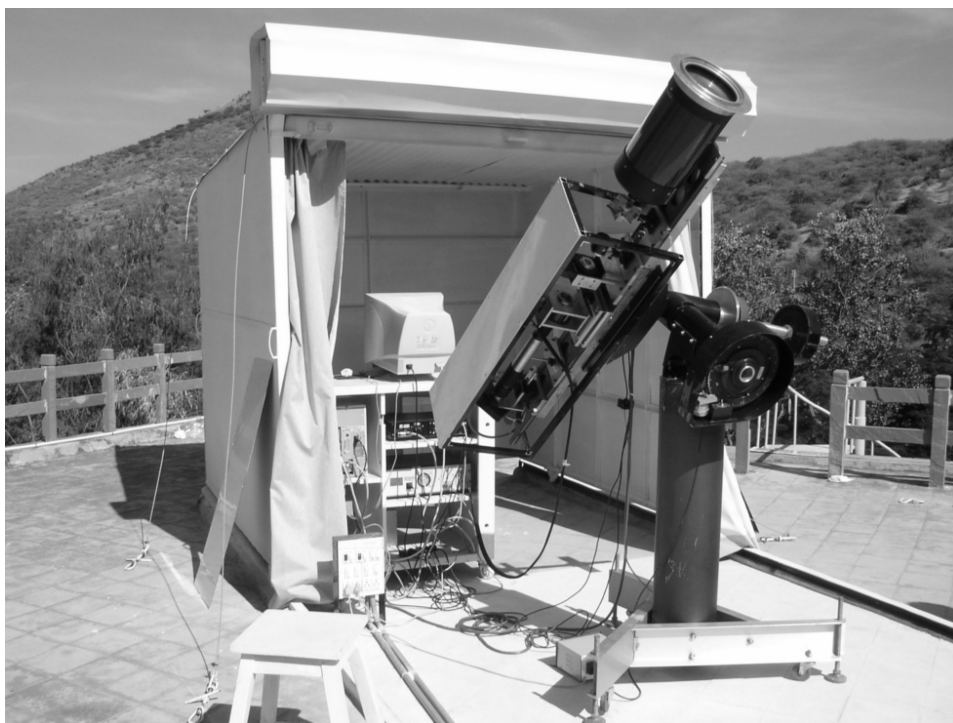


Figure 2. A view of the SVM Phase-I test set-up on the terrace of USO, Udaipur main campus building. The sliding dome is pushed backwards here and the SVM optics mounted on German Equatorial mount.

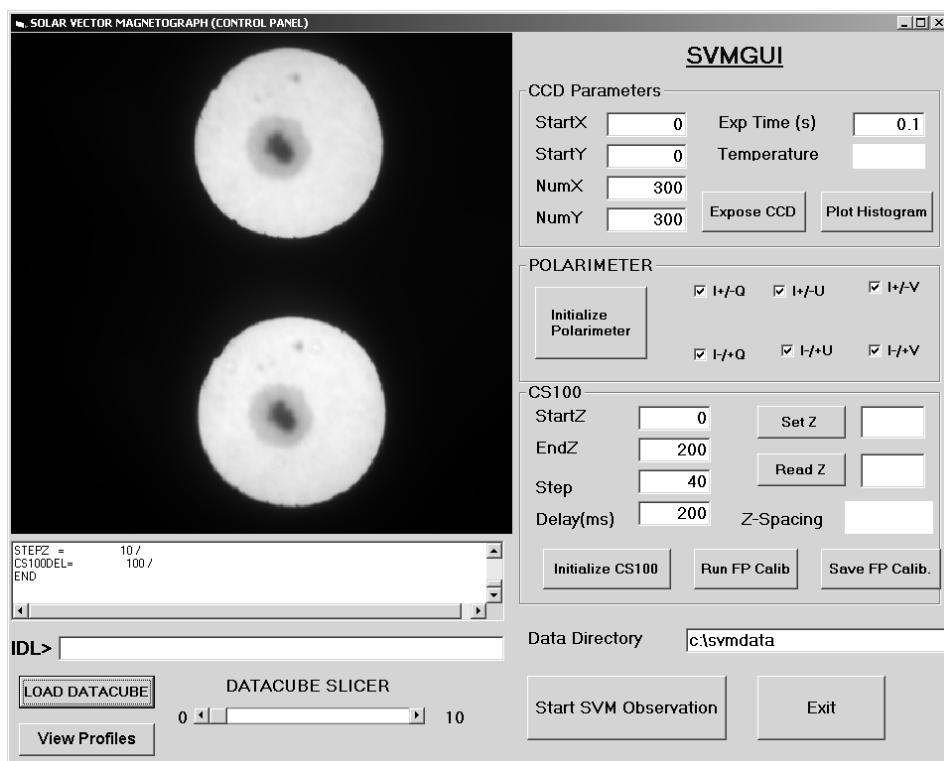


Figure 3. A snapshot of GUI for data acquisition and control of SVM instrument. The quicklook analysis of raw data can be done using IDL activeX on this interface.

2.5 Control software

The control software for the instrument has been developed in-house using Visual Basic on Microsoft Windows platform. The concept behind the software development has been to use ready-made software components like ActiveX drivers, COM modules, Dynamic Linked Libraries (DLLs), etc. from instrument vendors or third parties and integrate them together to be controlled under a single application. The CCD camera is controlled using ActiveX drivers supplied by the manufacturer. Similarly, the polarimeter rotary stages are controlled using COM modules from GPIB card vendor. The RS232 commands are sent using MSCOMM COM module. Hence, all the modules are accessed using one application. The application has a Graphical User Interface (GUI) which is used for passing parameters to the program. The images from the CCD camera are written in FITS (Flexible Image Transport System) format as three-dimensional data cubes and are visualised using ActiveX component of Interactive Data Language (IDL) data analysis software. A snapshot of the data acquisition and control software is given in Fig. 3.

3. Preliminary results

The first-light Stokes images of a sunspot from the instrument are shown in Fig. 4. These images show the Stokes signal in blue wing of the Fe I 630.2 nm line. Such

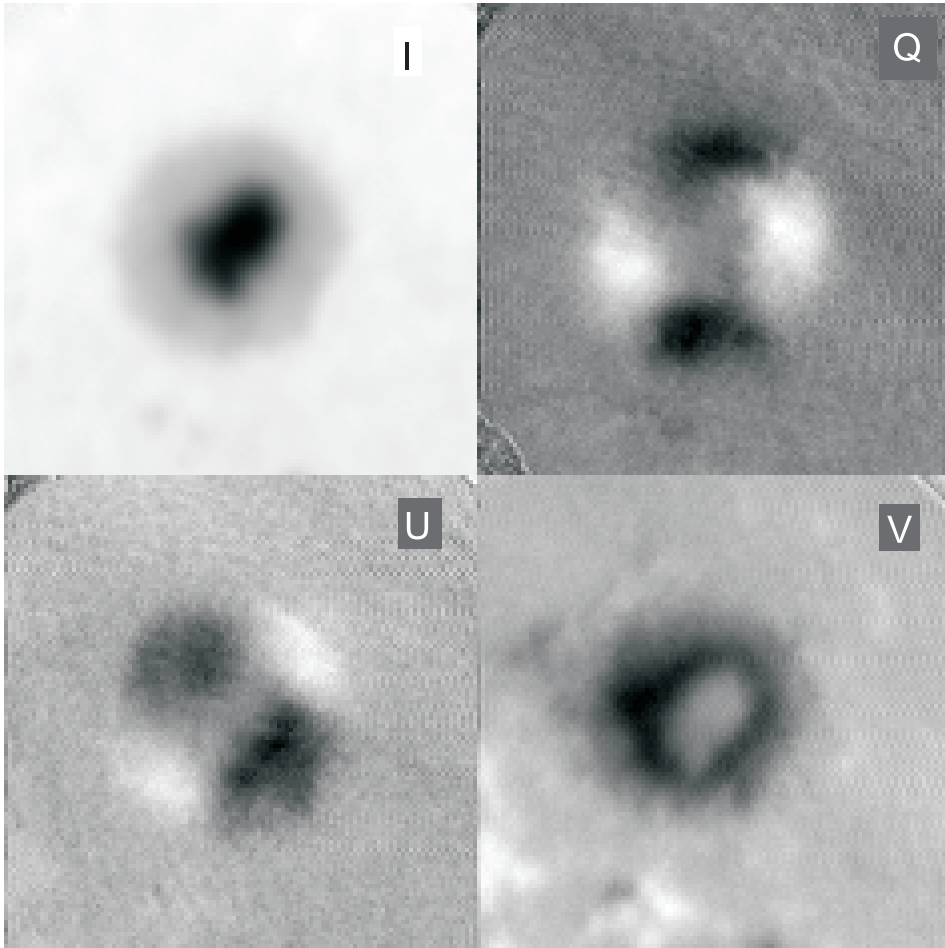


Figure 4. The first light observations represented here as Stokes I, Q, U and V images of a sunspot on the disk center. These images are obtained in the blue wing of Fe I 6302 line.

images are recorded all along the spectral line at discrete wavelength steps, the images are stacked to form a data-cube. The data-cube stores the Stokes profiles at all observed locations in the field-of-view. The data observed in this way is nearly free from seeing induced spurious polarization as these images are observed simultaneously. In principle the magnetic field vector can be deduced using the polarized images at only one wavelength position. This method relies on the calibration model to convert measured polarization to magnetic field. With measurements at only one wavelength position there are ambiguities, which are removed by measuring polarization at two or more wavelength positions in blue and red wings of the line (Hagyard & Kineke 1995). This method has been used successfully at various observatories around the world (Hagyard *et al.* 1982; Wang *et al.* 1998; Zhang H 2000; Sakurai *et al.* 1995) and is very suitable for studying magnetic field evolution in relation to solar flares. However, it is generally accepted that more accurate vector magnetic fields are derived by measuring full polarization spectral profiles. In this case, the observed profiles are inverted by

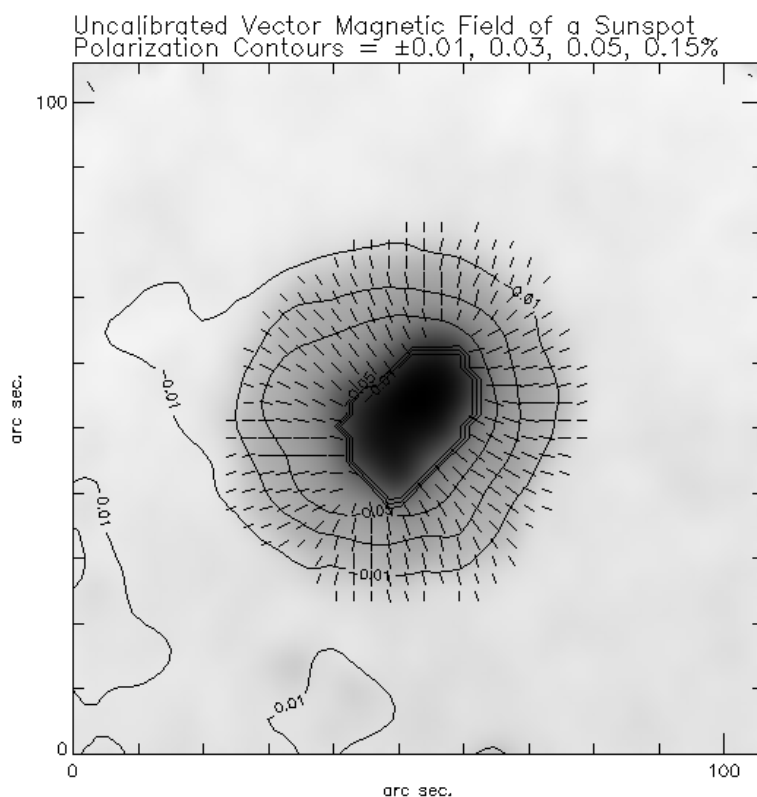


Figure 5. The uncalibrated vector magnetic field maps of a sunspot are shown using Stokes images of Fig. 4. The contours show the circular polarization degree, which is representative of line-of-sight magnetic field strength. The direction of the arrows represent the direction of transverse vector field.

non-linear least-squares fitting procedures, where a model atmosphere which yields observed profiles is searched. This approach is planned to be followed for the analysis of the SVM-I observations. Here, in a very preliminary analysis we present the uncalibrated vector field map in Fig. 5. The purpose is to check transverse field directions for their consistency with the configuration typically seen in sunspots. The observations are being further analyzed currently in order to determine the accuracies and sensitivity of the measurements, which will be published in a separate paper.

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Research Facilities for Solar Astronomy at ARIES

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Abstract. The solar observational facilities at ARIES (erstwhile U.P. State Observatory, UPSO), Nainital, began in the sixties with the acquisition of two moderate sized (25 cm, f/66 off-axis Skew Cassegrain and 15 cm, f/15 refractor) telescopes. Both these systems receive sunlight through a 45 cm and 25 cm coelostat respectively. The backend instruments to these systems comprised of a single pass grating spectrograph for spectroscopic study of the Sun and a Bernhard–Halle $H\alpha$ filter, coupled with a Robot recorder camera for solar patrolling in $H\alpha$ respectively. With the advancement in solar observing techniques with high temporal and spatial resolution in $H\alpha$ and other wavelengths, it became inevitable to acquire sophisticated instrumentation for data acquisition. In view of that, the above facilities were upgraded, owing to which the conventional photographic techniques were replaced by the CCD camera systems attached with two 15 cm, f/15 Coude refractor telescopes. These CCD systems include the Peltier cooled CCD camera and photometrics PXL high speed modular CCD camera which provide high temporal and spatial resolution of ~ 25 ms and ~ 1.3 arcsec respectively.

Key words. Sun: telescope—coelostat—solar activity—CCD—V++.

1. Introduction

Over the years the principal research interests in solar astronomy at the observatory have been focused primarily on studies related to molecules in photosphere, sunspots and faculae on the Sun and solar activities. These studies were mainly related to solar spectroscopy using single pass grating spectrograph and observations of solar activities in $H\alpha$ through a Bernhard–Halle $H\alpha$ filter. Various interesting results were obtained by using these facilities, few of them were reported by Sinha *et al.* (1979); Pande *et al.* (1980); Bondal *et al.* (1982); Pant (1984).

With passing time it was realized that the $H\alpha$ observations and associated X-ray and radio emission measurements, with high temporal and spatial resolution, were crucial for evaluating the energy release processes in the solar flares (Kiplinger *et al.* 1988). In order to achieve these objectives the CCD technology was introduced for the solar observations, with a time resolution of ~ 25 ms to 2 s and spatial resolution ~ 1.3 to 2 arcsec (Verma *et al.* 1996). Multi-wavelength studies of solar flares and associated radio bursts, CMEs, coronal holes, etc. in recent years have yielded some

interesting results about the solar activities (Anita Joshi *et al.* 2003; Uddin *et al.* 2004; Bhuwan Joshi *et al.* 2005). In the present paper we give a brief account of the telescopes, the backend instruments and methodology of data analysis in view of the solar observing facilities at ARIES, Nainital.

2. Chronology of solar observing facilities and methodology

The solar observational facilities at ARIES (erstwhile U.P. State Observatory, UPSO), Nainital were established in 1966 (Pande 1969). Then a 25-cm f/66 off-axis Skew Cassegrain telescope was acquired from Cox Hargreaves UK, attached with an observatory-made single pass grating horizontal solar spectrograph. The telescope forms an 18-cm solar image on the slit of the spectrograph. Except the grating, almost the whole of the spectrograph and the 46-cm coelostat (Fig. 1), were fabricated at optics and mechanical workshops of the observatory. The spectrograph was capable of providing a dispersion of $1.2 \text{ \AA/mm}$ in the first order at 2.5 micron. In addition to that, another system was acquired, consisting of a 25-cm coelostat and a 15-cm f/15 objective lens, forming an image of $\sim 22 \text{ mm}$ at the focal plane. The backend instrument for this telescope was a Bernhard–Halle $H\alpha$ filter ($0.5/0.7 \text{ \AA}$ pass band) coupled with a Robot Recorder 35-mm movie camera attached with a sequential timer, capable of providing a time resolution of ~ 0.15 second. The filter can be tuned to $\pm 1.0 \text{ \AA}$ of central wavelength in steps of $0.1 \text{ \AA}$.

These units were operational till 1988, when two Coude refractors (15 cm, f/15), from Carl Zeiss, Jena were acquired in 1988 and 1992. In the mean time a few narrow pass band DayStar filters, central wavelength at $H\alpha$ ($6562.8 \text{ \AA}$ pass band $0.5 \text{ \AA}$), Ca II K ($3933.7 \text{ \AA}$ pass band $1.2 \text{ \AA}$) and CN ($3883 \text{ \AA}$ pass band $1.0 \text{ \AA}$), were also acquired. The Coude refractor forms about 21 mm diameter image of the Sun, which is enlarged to 42 mm using a Barlow lens when the Bernhard–Halle $H\alpha$ filter is used. One of these telescopes has been put atop a 30-feet tower (Fig. 2). The filtergrams can be obtained on the Kodak Tech pan 2415 film with the help of 35 mm DN22 Robot camera, capable of recording 4 frames/second. With the prevalent solar observing techniques,



Figure 1. The 45-cm coelostat for the 25-cm, f/66 off-axis Skew Cassegrain telescope at ARIES.



Figure 2. The 15-cm Coude refractor telescope (inset box) put atop a 30 feet tower (left), to the right is the Coude refractor telescope for observations of solar activities.

with high temporal and spatial resolution, the aforesaid facilities were drastically improved in 1993 by adopting the video and CCD technologies in the solar observing facilities. The CCD technology was introduced for achieving the following scientific objectives:

- To record the solar activities (solar flares and prominences) with high temporal and spatial resolutions of ~ 25 ms and ~ 1.3 arcsec respectively.
- To compare the $H\alpha$ emitting regions during solar activities with those of the simultaneously observed microwaves and HXR solar emission data at high temporal and spatial resolution for understanding the flare triggering mechanism.

The entire upgraded system consists of a CCD chip, its camera controllers, a Sun Spark station-20 computer, a GPS clock for accurate timing, etc. The various CCD camera systems for solar observations have been described as under:

2.1 *Peltier cooled CCD camera system*

This CCD imaging system has EEV 385×578 pixel CCD chip (type P86231) with a pixel size of 22×22 microns. Both full frame (8.47×12.7 mm $\sim 6.3 \times 9.5$ arcmin on the Sun with Bernhald–Halle $H\alpha$ filter) readout and frame transfer modes are possible. The CCD is controlled through a PC and the data acquisition is done using ATI software.

2.2 *Photometrics PXL high speed modular CCD camera system*

This system has two camera heads attached with entirely two different cooling systems. One of them can be used at a given time for observing the Sun.

2.2.1 *Fibre bound CCD camera system*

This system has thermoelectrically cooled FTS cold probe camera head (ambient temperature -42°C) and TK 1024, Class I, CCD chip (pixel size 24×24 microns). Each pixel of the chip is connected to a window by fibre optics in such a way that one pixel receives light from a 2.45 times larger area at the window. Thus the camera head window acts as 58.8-micron sized pixel with a spatial resolution of approximately 5 arcsec. The camera controller with a 12-bit digitizer has variable readout rates from 0.5 to 1 mega pixel per second. The system is controlled by a Sun Spark 20, 50 MHz computer having 7.3 GB disk capacity and 14 GB cartridge tape drive (Verma *et al.* 1996). A GPS clock connected to the computer records time with an accuracy of 1 ms in the header of solar image, using a Sun interface software available with the above system. An electro-mechanical shutter, through an interface card fabricated at the observatory is used for computer-controlled exposures at the chosen instants and at the desired time intervals.

2.2.2 *High speed CCD camera system for flare observations*

The camera head of this system has EEV37 CCD chip (512×512 pixels, pixel size 15×15 microns) with a facility of frame transfer mask. The same camera controller, Sun computer software, etc. are used to obtain $H\alpha$ images of the solar flares. Liquid

circulation unit is used for cooling the camera head up to -25°C . The solar images having 100×100 pixels area, can be obtained with a time resolution of about 25 milliseconds or at the rate of 40 frames per second. Full frame images can be obtained at the rate of 5 frames per second (Gaur *et al.* 1998). The system is capable of providing a spatial resolution of ~ 1.3 arcsec which is extremely good for the solar flare observations. The data recorded by these CCDs are analyzed using IRAF and IDL software. We also have a COHO camera, connected to GPS clock for accurate time display, which is used for video recording of the full disk $H\alpha$ and white light images of the Sun. More than 20 research papers have been published based on the observations taken with these facilities.

Due to some problem in the SCSI controller of the photometrics PXL CCD camera system, the repairing and upgradation work of existing hardware and data acquisition software of photometrics PXL camera systems have recently been carried out at ARIES. The SCSI controller was replaced by adding AIA and PCI interface cards in the camera controller and PC respectively, facilitating the direct communication of the controller with the computer. This interface is also compatible with any of the windows operating systems. This modified system consists of P4, 3.2 GHz system with windows XP and having 1 GB RAM and 160 GB HDD. The program module has been made in Visual Pascal and Microsoft Visual C++ and this module is run in V++ software (by Digital Optics). The images are stored in FIT format. Therefore images can either be analyzed using V++ itself or any other software.

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Data Acquisition, Control, Communication and Computation System of Solar X-ray Spectrometer (SOXS) Mission

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Abstract. The Solar X-ray Spectrometer (SOXS) mission onboard GSAT-2 Indian Spacecraft was launched on 08 May 2003 using GSLV-D2 rocket by Indian Space Research Organization (ISRO). SOXS aims to study solar flares, which are the most violent and energetic phenomena in the solar system, in the energy range of 4–56 keV with high spectral and temporal resolution. SOXS employs state-of-the-art semiconductor devices, viz., Si-Pin and CZT detectors to achieve sub-keV energy resolution requirements. In this paper, we present an overview of data acquisition, control, communication and computation of low energy payload of the SOXS mission.

Key words. Data acquisition—X-ray spectrometer—instrumentation.

1. Introduction

Solar flares are explosive phenomena and they release gigantic energy in the solar system and thereby affect the Earth's environment and interplanetary space. Thus understanding solar flares is of the highest priority. The Solar X-ray Spectrometer mission is a step in this direction, in order to understand the energy release and particle acceleration mechanisms. The science objectives and instrumentation of the SOXS mission are described by Jain *et al.* (2000a, 2000b, 2002a, 2002b, 2005). X-ray spectroscopy of solar flares demands complex and state-of-the-art instrumentation. The solar corona is very hot, up to 5 keV and therefore X-ray flux in general is high and particularly during flares it becomes very high, while in high energy band (>10 keV) it starves for photons. Thus the designing of X-ray spectroscopy instrumentation, its operation and control becomes very complicated. We optimized the system in the dynamic energy range of 4–56 keV with the application of two detectors namely Si and CZT in order to observe almost all classes of flares and with satisfactory performances of the detectors. In section 2, we present very briefly data acquisition, control and communication aspects, while in section 3, we conclude our discussion.

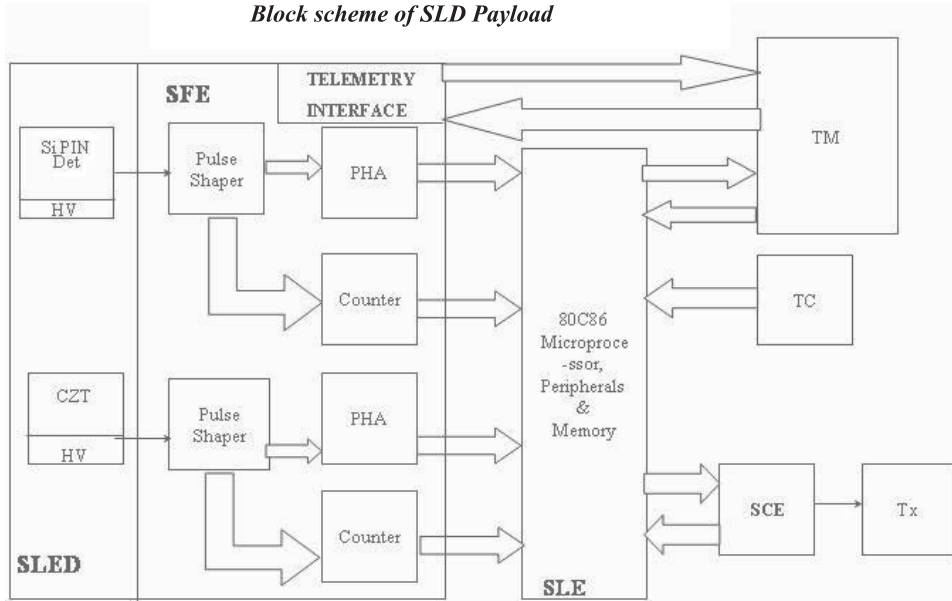


Figure 1. Block diagram shows SOXS low energy detector (SLD) subsystems. The signal flow from detector to telemetry packages is shown.

2. Instrumentation

In Fig. 1, block schematic of the low energy payload of SOXS mission is shown. The system consists of the following sub-systems:

- Low Energy Detectors (SLED) package consists of Si-Pin and CZT detectors.
- Front end Electronics (SFE) package consists of analog processing electronics.
- Low energy processing Electronics (SLE) package consists of data acquisition and control electronics.
- Common Electronics (SCE) package consists of power and communication electronics.
- Telemetry (TM), Tele command (TC) and Transmitter (TX) interface for spacecraft bus.

We discuss briefly the salient features in the following sections

2.1 Low energy detectors

This SLED package includes Si-Pin and CZT detectors with thermoelectric cooler, Charge Sensitive Pre-Amplifier (CSPA), high voltage dc-dc converters and heaters (Jain *et al.* 2002b). The detectors are mounted with collimators to restrict field of view to obtain better S/N ratio and to attenuate high energy charge particles. Each collimator has on-board calibration source for periodic calibration of the instrumentation. The details of the payload instrumentation is described in Jain *et al.* (2002a, 2002b, 2005). The SLED package is mounted on Sun Tracking mechanism (SSTM). SSTM controls SLED movement in RA and DEC axis with the help of single stepper motor which is

a very novel approach Viswanatha (2000). The SSTM is capable of positioning SLED with 0.1 deg. accuracy and can be controlled from ground station with the help of telecommands. Sun Sensors (SSS) are mounted on SSTM to get absolute position of the Sun. Two thermistors are also mounted on SSTM to monitor the absolute temperature of packages. The SLED is powered by SCE and the output signals of detectors and temperature sensors are provided to SFE for further processing.

2.2 Analog processing

This package includes linear pulse shaping amplifier, peak detectors, 8-bit ADC, energy window discriminator, telemetry interface for housekeeping (HK) parameters, temperature and corona auto-shut-off electronics (Jain *et al.* 2000b).

Pulse shaper and amplifier circuit consists of three stages of pulse integrating amplifiers. Analog pulse from SLED is shaped and amplified by this stage. Amplified pulse is compared with the Lower Level Discriminator LLD (4 keV) and Upper Level Discriminator ULD (25 keV for Si-PIN and 60 keV for CZT) signals. The comparators derive trigger output (Strobe pulse) if the input signal is within LLD–ULD range. This strobe pulse is the heart of SFE and used for event time generation.

Amplified signal of pulse shaper is split into two channels (1) high spectral resolution (82 eV/ch for Si-Pin and 212 eV/ch for CZT) ADC channel and (2) high temporal resolution (60 kps) window channel. Energy of pulse is classified into totally nine ranges to use them for flare detection logic.

Package health is monitored by 1 kHz telemetry. 16 parameters are monitored every 16 sec through LBT. The parameters monitored are detector temperatures, SFE dc-references and salient threshold voltages. Total events recorded in both detectors are also monitored by telemetry.

Temperature auto-shut-off circuit switches off the detector bias voltage when the detector temperature falls out of the operating temperature range of -25°C to $+5^{\circ}\text{C}$. This will protect the detector from thermal breakdown. Corona auto-shut-off circuit switches off the detector bias when the detector bias current increases 50% above the normal value.

2.3 Data acquisition and control

This package consists of 16-bit window counters, ADC interface, real-time pulse height analysis (PHA) logic, 5 MB on-board memory to store flare and HK data, On-Board Timer (OBT) and telemetry and telecommand interfaces (Umapathy *et al.* 2003).

The PHA generates real-time histogram from ADC data and stores it in PHA memory. The PHA memory is flushed every sampling time to on-board flare memory. The timing logic generates all required time in such a way that each event is processed within 20 μsec . The nine window pulses are gated to nine 16-bit counters. These counters are sampled continuously to decide flare on-set and stored every sampling time to on-board flare memory. Two out of nine counters are used and compared with up-linkable thresholds to decide about quiet or flare phase. SFE HK parameters are routed to telemetry via processing electronics.

The SOXS processing electronics has a total of 5 MB on-board memory divided into two banks each of 2.5 MB. The instrumentation provides ‘a continuous data recording and playback’ mechanism. When one bank is busy with acquiring data the other is

flushing data to telemetry at 8 kHz rate. The banks are switched every 29.8 minutes for data acquisition.

On-board 16-bit CPU controls all operations of processing electronics. The SOXS current operating mode is programmable via telecommand. Operating modes are:

- Flare mode (search, quiet and flare integrated),
- Survey/background mode (100 ms and 1 s),
- Memory and electronics check-out mode,
- Readout mode.

Flare mode is the normal operating mode. In this mode, search is going on for flare onset. Quiet phase PHA and counters data are stored every 3 s and 1 s respectively. If flare is detected then data stores every 100 ms. The quiet and flare phases data are packaged into 700 bytes packet.

The packet consists of:

- Header (packet ID, time tag, flare status) information,
- Pulse height analysis (real-time spectrum) information,
- Energy window counters (temporal) information,
- Padding information.

Survey mode is used to study the level of background noise of quiet sun. This helps to estimate correct thresholds for flare detection.

Memory check-out and read out modes are usable for on-board diagnosis.

2.4 Data communication

This package provides a common interface to all SOXS packages with spacecraft bus. It minimizes the chance of damage of mainframe bus because of anomaly in packages. This package consists of low voltage dc/dc converters, high current ART converters, power electronics, telemetry/telecommand interfaces, base band data coder, convolution encoder, pre-modulation filter, sun sensor and SOXS mechanism drive electronics (Sharma *et al.* 2000).

Two sets of RMUs and digital MUX provide common telemetry interfaces to SLED, SFE and SLE packages. Convolution encoding is an error correction technique used in data communication. This encoder can be bypassed by telecommanding in case error correction is not required. All high bit rate (HBT) and low bit rate (LBT) telemetry interfaces are conditioned in this package for final transmission to spacecraft telemetry. Telecommand received by spacecraft will be first received by this package and decoded and DEMUXed for respective packages.

2.5 Data computation

The 700-byte packets are continuously uploaded to Master Control Facility (MCF) at Hassan, India. The packet is again combined with other monitoring parameters at MCF and 1248-byte packets are formed.

The final packet consists of:

- Frame ID (sync. bytes, data ID, frame counter),
- Data packet status,

- Derived and current PID values (parameters at onboard and current time),
- Frames Ground Reception Time (GRT),
- Dummy bytes.

The 1248 packet consists of housekeeping parameters and flare data.

The MCF computer is linked to PRL through space net. Space net is the dedicated 64 Kbps peer-to-peer network commissioned for space data interchange within ISRO units. The following computation is done on raw data.

- Extraction of health parameters,
- Instrumental correction,
- Validation of data w.r.t. health parameters,
- Averaging and construction of spectra,
- Gaussian/linearity curve fitting,
- Effective area calculation/energy calibration,
- Detector response fitting.

The IDL package installed on Linux enabled PCs are mainly used for off-line data computation.

3. Conclusion

The SOXS mission was successfully launched, and in-orbit tests were conducted successfully. On-board subsystem parameters were evaluated and found satisfactory as per requirements. ∇T of the detectors found lower than predicted from thermal design restricts observing duty cycle to 3.5 h only. Satellite grounding noise shift CZT spectrum by 4 keV which restricts dynamic range to 4–56 keV. Apart from these detectors, analog processing, data acquisition, control and communication subsystems performance were found very satisfactory.

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Observational Study of Solar Magnetic Active Phenomena

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Abstract. The electric current separated into two parts reflected the quantitative properties of heterogeneity and chirality of magnetic field, and defined them as the shear and twist components of current. We analyze the basic configuration and evolution of superactive region NOAA 6580-6619-6659.

It is found that the contribution of the twist component of current cannot be reflected in the normal analysis of the magnetic shear and gradient of the active regions. The observational evidence of kink magnetic ropes generated from the subatmosphere cannot be found completely in some super delta active regions.

Key words. Sun: activity—flares—magnetic fields.

1. Introduction

It is believed that the newly emerging magnetic flux of opposite polarities and the shear of transverse magnetic field near the magnetic neutral line in active regions are important parameters for the non-potential magnetic field (Hagyard *et al.* 1984). The non-potentiality relates to the existence of electric current and the trigger of solar flares (Lin and Gaizauskas 1987).

2. Magnetic shear, twist and relationship with electric current

The electric current reflects the non-potentiality of magnetic field in the solar atmosphere and can be inferred from the magnetic field in the form

$$\mathbf{J} = \frac{1}{\mu_0} (\nabla \times \mathbf{B}), \quad (1)$$

where $\mu_0 = 4\pi \times 10^{-3} \text{ GmA}^{-1}$.

The magnetic shear or twist normally reflects the existence of electric current and helicity in the active region. The electric current can be written in the form (Zhang 2001)

$$\mathbf{J} = \frac{1}{\mu_0} (\nabla B) \times \mathbf{b} + \frac{B}{\mu_0} \nabla \times \mathbf{b}, \quad (2)$$

where $\mathbf{B} = B\mathbf{b}$ and $\mathbf{b}$ is the unit vector along the direction of magnetic field and $B = \sqrt{B_x^2 + B_y^2 + B_z^2}$. This provides that the electric current in solar active regions relates to the properties of gradient and chirality of magnetic field. The first term in equation (2) connects with the heterogeneity and orientation of magnetic field, *i.e.*, the shear of vector magnetic field. The second term in equation (2) connects with the twist of unit magnetic lines of force and the intensity of field. We can artificially define that the first term is the shear current and the second term is the twist current in equation (2). According to equation (2), the vertical current can be inferred from

$$J_z = \frac{1}{\mu_0} \left(b_y \frac{\partial B}{\partial x} - b_x \frac{\partial B}{\partial y} \right) + \frac{B}{\mu_0} \left(\frac{\partial b_y}{\partial x} - \frac{\partial b_x}{\partial y} \right). \quad (3)$$

3. Delta active region NOAA 6580-6619-6659

3.1 Formation of large-scale non-potential magnetic field

Figure 1 shows the sunspots and corresponding vector magnetic field in this large delta active region NOAA 6580-6619-6659 observed at Huairou Solar Observing Station of National Astronomical Observatories of China in 1991 (Ai & Hu 1986). It is found that the sunspots of opposite polarities grew with the emergence of magnetic flux from the sub-atmosphere to form a bipolar region, and became a compact delta region gradually. The twist of transverse component of photospheric magnetic field increased with the development of active region. As the shear angle of magnetic field is defined by that between the connected line of large-scale magnetic poles of opposite polarities and the transverse field near the magnetic neutral line, one can find that the shear angle of transverse field increased from about 30° on April 14 to 44° on June 9. The flux of magnetic poles of negative polarity increased faster than positive one in active region NOAA 6580-6619-6659. The total magnetic flux in active region NOAA 6580-6619-6659 increased from 2.7×10^{21} Mx to 2.2×10^{22} Mx (maximum value of this active region) during the period between April 14 and June 9, 1991. The total flux increased more than 8 times.

3.2 The relationship between the magnetic shear, gradient and current

The magnetic shear is an important parameter to measure the non-potentiality of magnetic field in solar active regions (Wang *et al.* 1994), while the non-potential field can also be measured from the strong magnetic gradient of active regions, which is strongly correlated with active region CME productivity (Falconer 2001).

Figure 2 shows the distribution of magnetic shear in the active region 6659 on June 9, 1991. The shear is defined from the shear angle weighted by transverse magnetic field

$$\theta_T = B_T \cdot \cos^{-1} \frac{\mathbf{B}_T \cdot \mathbf{B}_{pT}}{B_T B_{pT}}, \quad (4)$$

where B_T and B_{pT} are observed transverse field and calculated one from the magnetic charges in the photosphere. The amplitude of the shear angle relates to that of the transverse field and the angle difference between the observed transverse field and

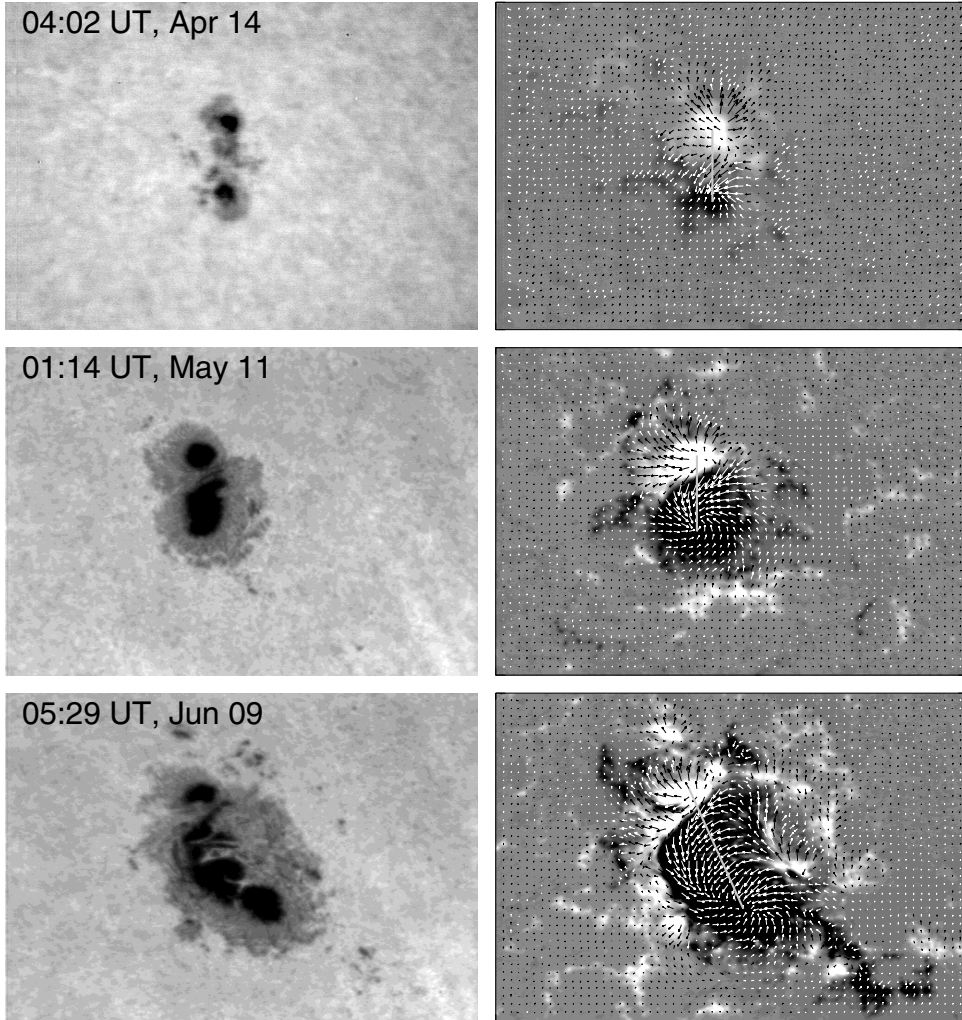


Figure 1. The white light images (left) and vector magnetograms (right) in active region NOAA 6580-6619-6659 in 1991. The white (black) areas indicate positive (negative) polarity. The horizontal arrows mark the transverse component of magnetic field. The straight lines connect the magnetic main poles of opposite polarities. North is at the top, west is at the right. The size of images is $5.23' \times 3.63'$.

inferred potential one. It is noticed that the definition of the shear angle of the transverse field in equation (4) is slightly different from the first term of equation (2), but the physical means are almost the same.

The gradient of photospheric longitudinal magnetic field in the active region can be inferred from

$$|\nabla(B_z)| = \sqrt{\left(\frac{\partial B_z}{\partial x}\right)^2 + \left(\frac{\partial B_z}{\partial y}\right)^2}. \quad (5)$$

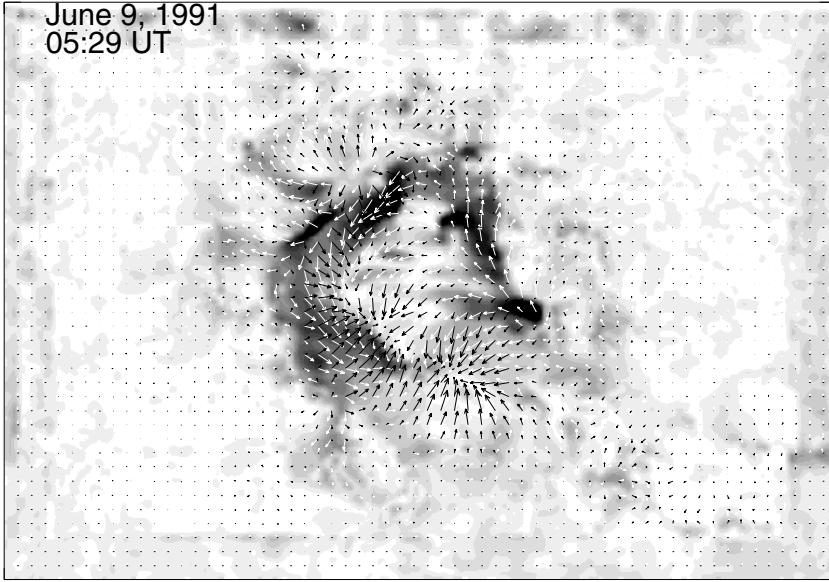


Figure 2. The distribution of magnetic shear in active region NOAA 6659 on June 9, 1991. The white arrows mark the observed transverse field and the black arrows show the transverse components inferred from the calculation of magnetic charges.

The distribution of the magnetic shear and corresponding magnetic gradient in active region NOAA 6580-6619-6659 inferred from photospheric vector magnetograms is shown in Fig. 3. The main contribution of magnetic shear in the active region comes from the deviation of the transverse field from the potential field inferred by magnetic charges in the photosphere, while the magnetic gradient comes from the non-uniformity of the longitudinal field.

By comparing equations (4–5) with equation (2), it is found that the basic information on the magnetic shear and gradient are contained in the shear term of the current. The definition of the angle of observed magnetic shear in equation (4) relative to the direction of transverse field inferred from the potential field is slightly different from the orientation of transverse field relative to ∇B , but it reflects the difference of the reference systems in analyzing the non-potentiality of magnetic field only. This means that the shear component of the current actually provides both magnetic shear and gradient of field in active regions. The directions of magnetic shear and gradient are lost in equations (4–5). Moreover, it is also noticed that the contribution of twist component of the current is not contained in the analysis of the magnetic shear of transverse field and the gradient of longitudinal field in the active regions. This means that the electric current is a relative complete quantity than the magnetic shear and gradient in the study of the non-potential field. The magnetic shear and gradient are not enough to describe the magnetic non-potentiality completely.

3.3 Comparison between twist and shear components in the current

Figure 4 shows the distribution of shear and twist components in the electric currents in active region NOAA 6580-6619-6659 inferred from photospheric vector

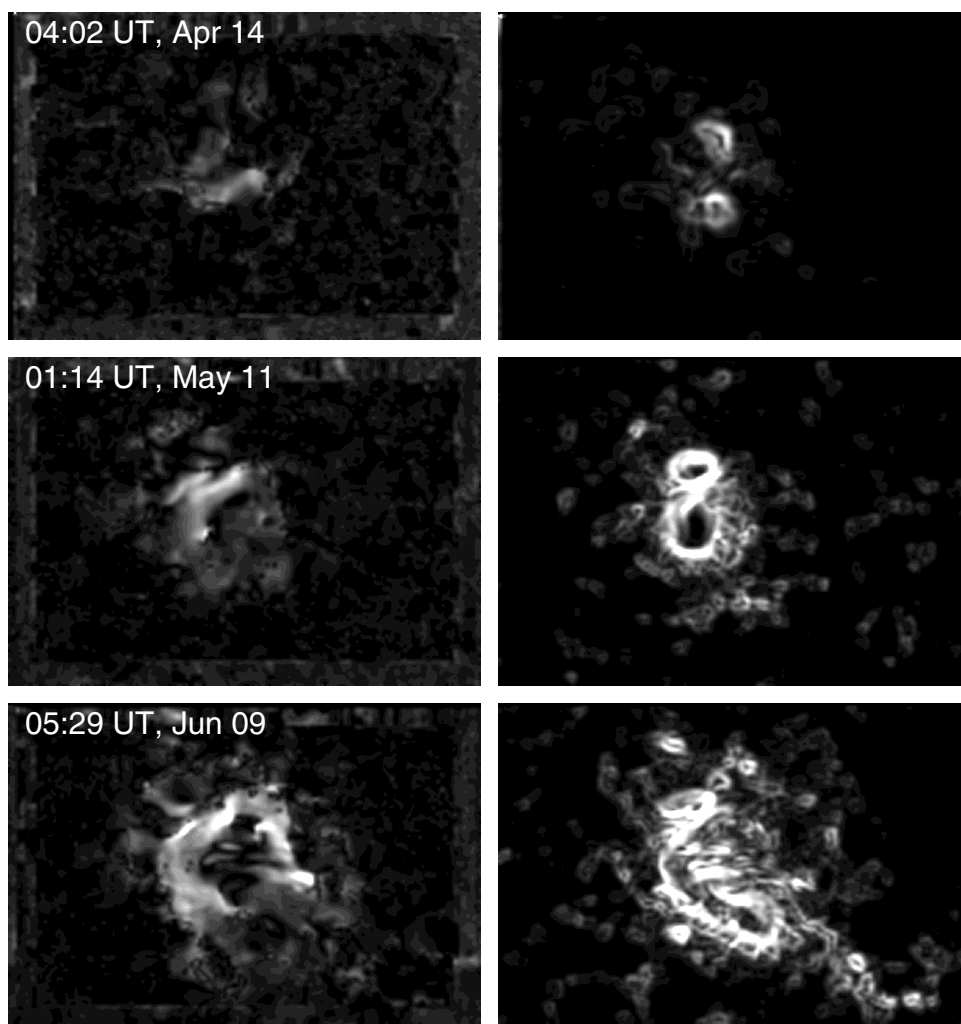


Figure 3. The magnetic shear (left) and gradient (right) in active region NOAA 6580-6619-6659 in 1991 inferred from photospheric vector magnetograms in Fig. 1.

magnetograms. The shear and twist components of the current contain more information on the intrinsic property of non-potentiality of magnetic field in the active regions than the total current. We can find that the topological distribution of both components of the current and their basic relationship relates with the development of active regions, *i.e.*, the topology of electric current in the photosphere obviously connects to the amount of magnetic flux in the active region. The shear current probably consists of an observational evidence of that current flows near the interface of the magnetic poles of opposite polarity and the reason why the inversion line of the vertical electric current obviously crosses the magnetic inversion line in some active regions (Zhang 1997). The mean ratio of statistical standard errors between the shear and twist current in NOAA 6580-6619-6659 on April 14–16 is 0.80, on May 9–13 is 0.54 and on June 8–9 is 0.5. This means that the increment of shear component of the current with

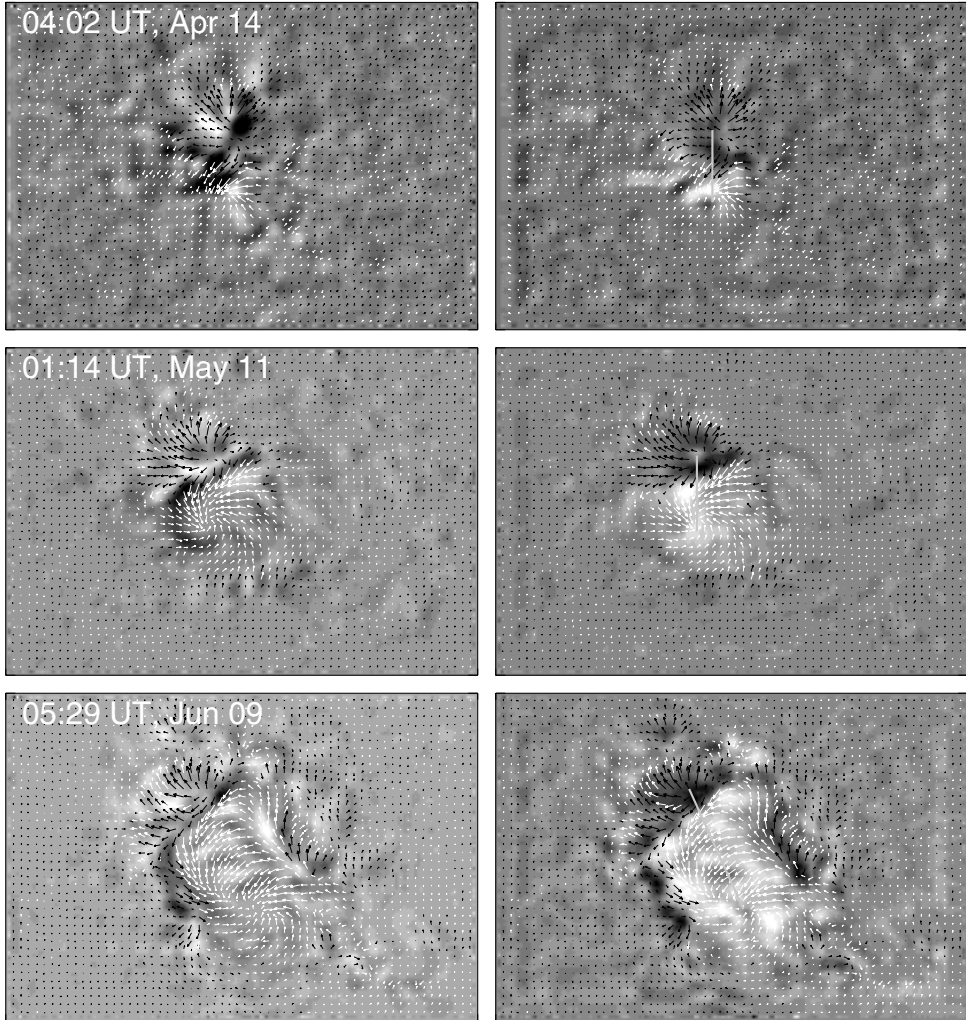


Figure 4. The vertical components of shear (left) and twist (right) component of current in active region NOAA 6580-6619-6659 inferred from the corresponding vector magnetograms of Fig. 1. The white (black) indicates the up (down) flow. The horizontal arrows indicate the transverse components of magnetic field. The straight lines connect the magnetic main poles of opposite polarities.

the development of active region NOAA 6580-6619-6659 is slower than twist one. The twist component of the current shows a dominant tendency of anti-correlation with longitudinal magnetic field, while shear component of the current does not. We can find a basic conclusion that the twist component of the current is the dominant one for the total current in the active region. A similar case has been demonstrated by Leka and Barnes (2003a, b) also. The significant contribution of the shear component of the current near the magnetic neutral line in the early stage of the active region probably reflects that the portion of non-force free field is more notable than the late stage (Georgoulis & Labonte 2004), as one believes that the shear and twist

component of the vertical currents provides the different basic information of the total current (Zhang 2001; Leka & Barnes 2003a, b) and the shear component of the current is not cancelled by the twist one completely.

3.4 Twist and writhe of magnetic ropes

The magnetic twist and writhe are two basic quantities reflecting the properties of solar helical magnetic loops, which probably generated in the solar sub-atmosphere. Linton *et al.* (1999) proposed that a concentrated kink would first emerge highly tilted, rotate during its subsequent emergence while remaining compact and develop strong shear along its magnetic neutral line. The main characteristic of the kink instability is that the flux loop axis is distorted having the same handedness as the twist. This mechanism necessarily implies that the sign of twist and writhe should be the same (Linton *et al.* 1999; Lopez Fuentes *et al.* 2003). While, it is found that the mean twist and writhe (relates to the tilt angle of active regions) of magnetic field in most bipolar active regions show opposite signs (Tian *et al.* 2001), the opposite statistical result is obtained by Holder *et al.* (2004).

In the northern hemisphere, the right-handed writhe of magnetic loops caused by Coriolis force probably implies that the magnetic field in the photosphere should twist clockwise (contrary-clockwise) in the positive (negative) poles due to the conservation of magnetic helicity. The magnetic field in active region NOAA 6580-6619-6659 twisted in left-handedness in the magnetograms is shown in Fig. 1. The shearing motion of magnetic weighting center of positive (negative) polarity towards east (west) direction in Fig. 1 implies that the writhe of magnetic loops is right-handedness (Lopez Fuentes *et al.* 2003). By comparing the mean sign of the handedness of twist magnetic field (magnetic helicity) and the evolution of active region (the relative moving direction of the magnetic main poles), one can find that active region NOAA 6580-6619-6659 belongs to the regular one, i.e., it is the opposite signs between twist and writhe of magnetic field.

As one analyzes the evolution of active region NOAA 6580-6619-6659 by the Daily Activity Solar Maps in Solar-Geophysical Data, it is found that the tilt angle of the active region was about 68° on April 14, 52° on May 11 and 46° on June 9, while the corresponding force free factor α_{best} was -3.3 , -4.2 and -4.5 (unit is 10^{-8} m^{-1}) for the areas that the transverse field is larger than 200G in vector magnetograms and corresponding mean current helicity density $\overline{h_{\text{CZ}}}$ was -0.22 , -4.8 and -5.2 (unit is $10^{-4} \text{ G}^2 \text{ m}^{-1}$). The similar tendency of tilt angles of the active region can also be found in the Kitt Peak synoptic magnetic maps. It is found that the reference tilt angles were about 86° , 54° , 50° , 22° , 28° and 14° in 1841, 1842, . . . , 1846 Carrington rotation respectively, even if in 1844, 1845 and 1846 Carrington rotation the active region became the enhanced magnetic networks. The tilt angle of the active region decreased and the twisted magnetic field increased during the development of active region. This means that the mean writhe magnetic field does not relax with the twist one synchronously.

4. The formation of non-potentiality of the magnetic field

Figure 5 shows an example of $\text{H}\beta$ flares in the active region on May 10, 1991. The $\text{H}\beta$ flare ribbons occurred in both sides of the magnetic neutral line. A ribbon consisted

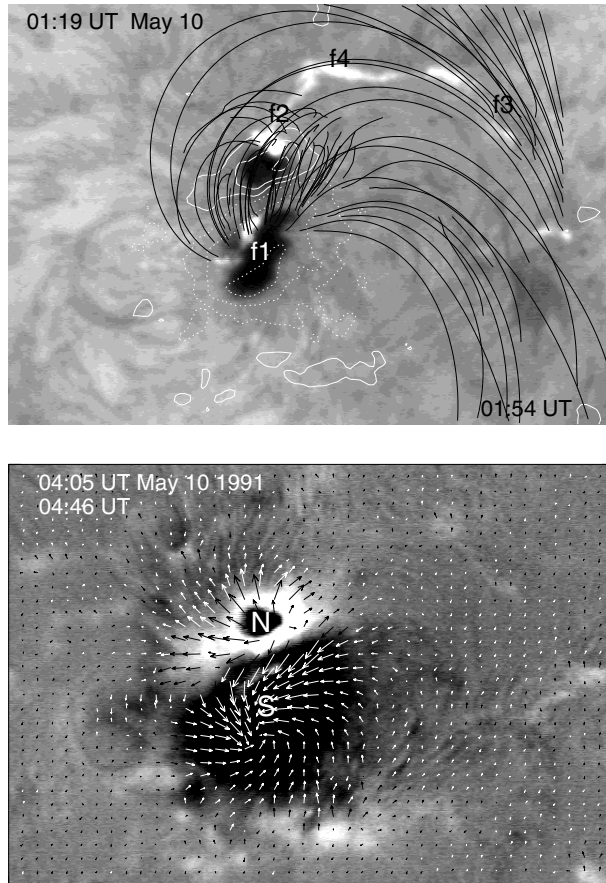


Figure 5. The $H\beta$ flare in active region NOAA 6619 on May 10, 1991 overlapped by the extrapolated magnetic lines of force (top). The solid (dashed) contours mark the positive (negative) polarities of the photospheric magnetic field. The chromospheric longitudinal magnetic field and photospheric transverse magnetic field (bottom). The white (black) marks the positive (negative) polarity.

of two cores marked by $f1$ in the umbra of negative polarity and another one marked by $f2$, $f3$ and $f4$, where $f2$ was located in the umbra and $f3$ and $f4$ extended into the enhanced network of positive polarity. It is found that the extrapolated magnetic lines of force connect both the ribbons of opposite polarities under the approximation of linear force free field. The α factor is -4.0 (unit is 10^{-8}m^{-1}) for the extrapolated magnetic lines of force from the magnetogram on May 10, 1991. In the observational photospheric vector magnetogram, the highly sheared transverse component formed in the magnetic neutral line between the flare ribbons with opposite magnetic polarities marked by N and S in Fig. 5. Comparing the relationship between the $H\beta$ magnetogram and photospheric transverse field, it is found that the chromospheric magnetic features around the magnetic main poles N and S show a tendency that these features are almost parallel to the direction of the photospheric transverse field. It reflects a basic evidence on the consistence between the distribution of chromospheric magnetic fibrils and direction of observational transverse field resolved 180° – ambiguity near the center of

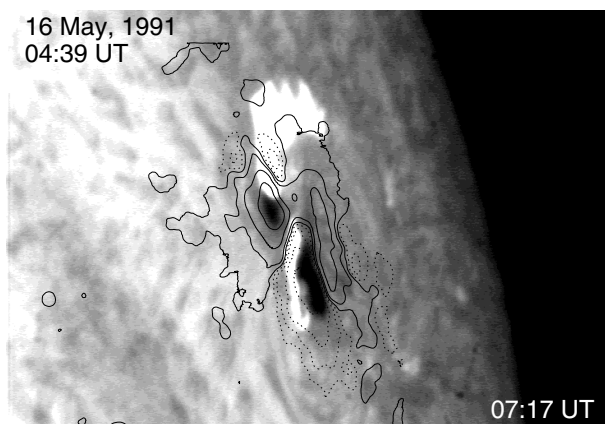


Figure 6. The $H\beta$ flare in active region NOAA 6619 on May 16, 1991 overlapped by the photospheric longitudinal magnetic field. The solid (dashed) contours mark the positive (negative) polarities.

solar disk. The reversal structure in the umbra N in the chromospheric magnetograms probably is caused by the disturbance of photospheric blended lines in the wing of $H\beta$ line (Zhang 1996). Figure 6 shows a limb $H\beta$ flare in the active region on May 16, 1991 in which flare ribbons show the similar pattern to that of May 10, even if the projective effect of the longitudinal magnetic field influences the adjustment of the relationship between the position of flare ribbons and magnetic polarities. As the comparison with development of transverse magnetic field near magnetic neutral line in the active region in Fig. 1, it is found that the shear of magnetic field increased after flares on May 10, 16 in Figs. 5 and 6. The similar observational evidence was also demonstrated by Wang *et al.* (1994) and Zhang *et al.* (1994). It means that the increase of highly sheared magnetic field near the magnetic neutral line and the interaction with the existed field are basic processes in the development of active region NOAA 6580-6619-6659, even if a series of powerful flares occurred in it. The shear increase after powerful flares contradicts the kink model of twisted magnetic ropes. It is hard to infer that high intense kinked magnetic ropes formed after the reconnection of magnetic field related to the powerful flares.

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Development of Solar Scintillometer

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Abstract. The index of scintillation measurement is a good parameter to compare different sites for image quality or ‘seeing’. We have developed a scintillometer, which is deployed on the high resolution SPAR telescope in the island site of Udaipur Solar Observatory, for the site characterization to specify the proposed MAST (Multi Application Solar Telescope).

The scintillometer consists of a miniature telescope, termed as micro telescope (4 mm aperture, 15 mm focal length) mounted on a drive which tracks the Sun continuously, associated amplifiers and a data acquisition system. A photodiode is used as the detector. The telescope along with detector was obtained from National Solar Observatory (NSO), and is similar to the one used for Advanced Technology Solar Telescope (ATST) site survey. At USO we developed the amplifier and data acquisition system for the scintillometer. A 24-bit analog to digital converter based system was designed, assembled, tested and used as the data acquisition system (DAS). In this paper, we discuss the instrumentation and present the initial results.

Key words. Sun: scintillation, scintillometer, seeing.

1. Introduction

Atmospheric turbulence causes refractive index inhomogeneities in the air. Beams passing through the turbulent air are constantly being refracted. The net result of multiple deflections along the ray path results in blurring and scintillation. It is found that scintillation in the solar irradiance is closely related to image quality. Scintillometer measures the intensity variation in the incident beam caused by atmospheric turbulence. By measuring the scintillation it is possible to evaluate seeing for the site (Seykora 1993; Beckers 1998).

2. The scintillometer

The scintillometer which we developed can be basically divided into three parts.

- micro telescope
- signal processing unit and
- a data acquisition system.

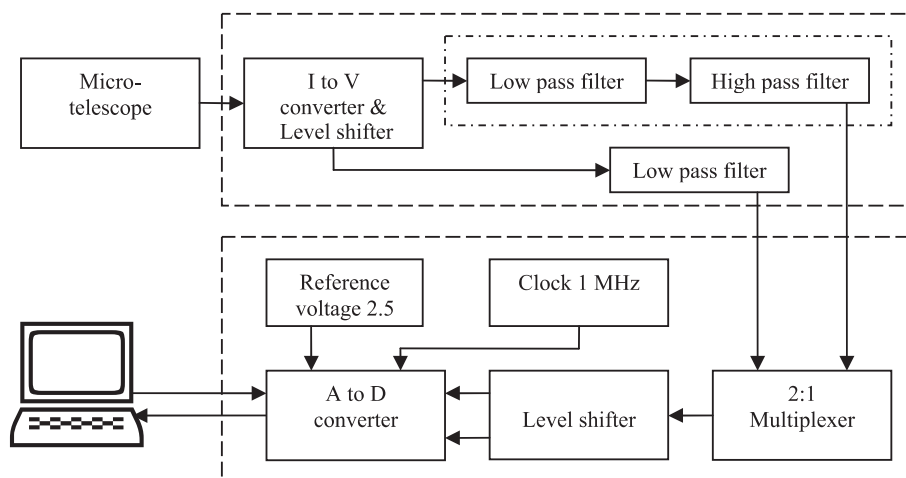


Figure 1. Schematic of scintillometer system.

The micro telescope collects the light from around 28 solar radii, including the Sun and feeds the collimated light into a photo-diode. The intensity of the collimated light consists of two components, a slowly varying (*dc*) component due to the diurnal variation of solar irradiance, and a superimposed rapidly changing (*ac*) component in Sun light intensity due to atmospheric effects, which is termed as scintillation. Since the amplitude of the *ac* component is relatively small, the measurement of scintillation involves extraction and amplification of the *ac* component from the total intensity signal. Two channels in the data acquisition system acquire both these components almost simultaneously with a sampling rate of 30 Hz. Figure 1 shows the schematic of the whole system and the complete circuit diagram is shown in Fig. A1. In the following sections, various components of the scintillometer are described in detail.

2.1 Micro telescope

The micro telescope is mounted on a tracking system. It consists of an objective lens of size 5 mm and effective focal length of 15 mm followed by two filters (blue and clear). A 2 mm field stop and a Fabry lens form an image of 4 mm diameter entrance aperture on a Silicon pin photodiode. The wavelength coverage is 220 nm centered around 510 nm. The pin photodiode output is connected to the signal processing unit using shielded cables and BNC connectors.

2.2 Signal processing unit

The signal collected by the photodiode is fed to a signal processing unit. This signal processing unit consists of a current-to-voltage converter, filter circuits to separate out *dc* and *ac* components, and amplifiers to boost the signal level. The circuit is based on discrete components which include op-amps (TL082, TL084) and other passive parts. A low-pass filter with a cut-off frequency of 0.3 Hz is used to retrieve the slowly varying *dc* component while a band-pass filter with on and off frequencies of 0.1 Hz and 700 Hz respectively, is employed to separate out the scintillation signal. The output voltage

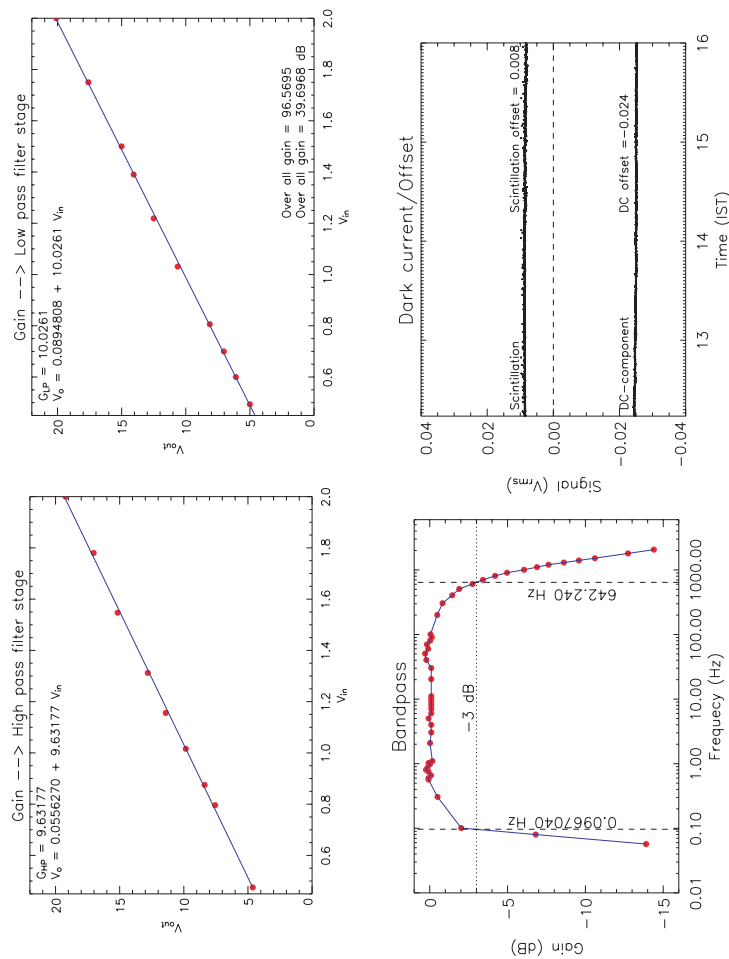


Figure 2. Calibration plots for signal processing unit and A/D converter. **Top:** Gain measurement for low-pass and high-pass filter stages. **Bottom:** (left) Frequency response of the band-pass filter, (right) dark current calibration for A/D converter.

from the *dc* part of the amplifier is adjusted such that, it does not exceed 8 volts during the entire day. The scintillation signal is amplified by a factor of 96 with respect to the *dc* signal, and for typical seeing conditions the value lies between ± 2 volts. Figure 2 shows the gain for the *dc* and *ac* amplifier sections and frequency response/band pass of the *ac* section.

2.3 Data acquisition system

A 24-bit A/D converter based (Burr-Brown, ADS1252) data acquisition system is developed for digitizing the data. A 2.5 volts reference and auxiliary level shifting circuit makes the A/D converter capable of digitizing input voltages up to ± 10 volts which is quite sufficient for our purpose. The A/D converter requires two separate clock signals CLK, SCLK for digitization and readout, respectively. A 1 MHz crystal oscillator circuit provides the CLK pulses, while readout clock pulses are generated by a computer and is applied through its parallel port to the A/D converter. The computer generates all other signals required for the A/D converter and also reads and stores the digitized data. The readout speed is limited to 70 Hz due to the slow response of the parallel port, and in our case it is carried out with a frequency of 30 Hz for each channel. A 2:1 multiplexer alternates the input signals (*dc* and *ac* components) to the A/D converter.

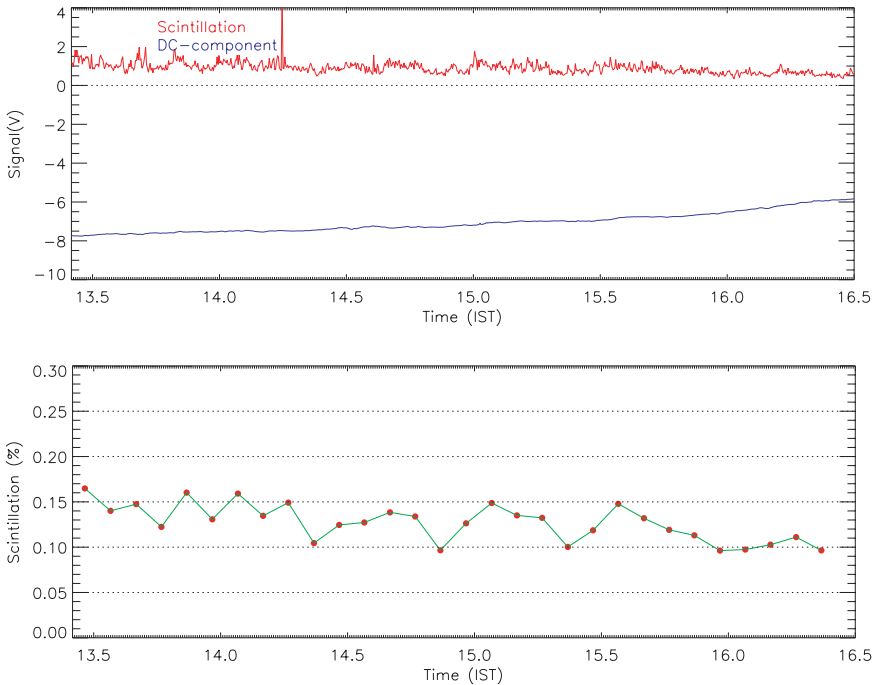


Figure 3. Scintillation recorded on March 31, 2005. **Top:** *rms* scintillation integrated over 10 seconds and the average *dc* component. **Bottom:** percentage scintillation averaged over 6 minutes.

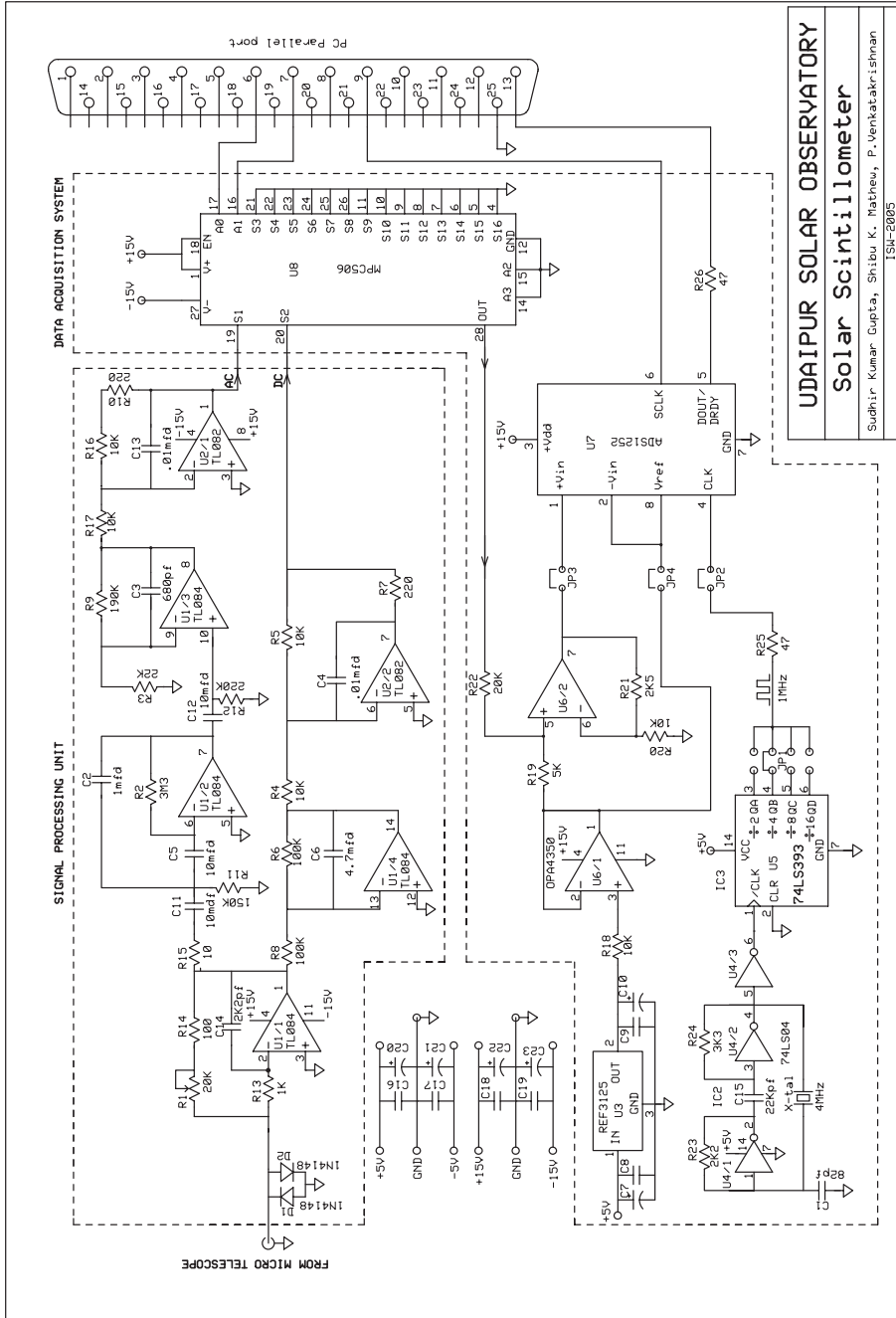


Figure A1. Scintillometer circuit including the signal processing unit and data acquisition system.

Programs are developed in C++ and IDL for operating the data acquisition system from the computer. The C++ program produces all required clock pulses and carries out the digitized data readout. A Graphical User Interface (GUI) is developed in IDL which collects all user inputs such as readout frequency, integration time, etc. and passes on these parameters to the C++ program. The free parameters which can be adjusted are, the start and end time for data recording, channel selection, readout frequency and integration time. The GUI creates a data file for each integration segment and also calculates and stores the rms value of scintillation for the entire observing period.

3. Preliminary results

Figure 3 shows the preliminary result of scintillation measurements. These measurements were carried out on March 31, 2005 from 13:30 to 16:30 IST. The top panel shows *rms* scintillation integrated for 10 seconds and the *dc* component. In the bottom panel we display the percentage scintillation averaged over 6 minutes for the same time period. The percentage scintillation is computed by taking the ratio of *ac* and *dc* components after normalizing the *ac* component with the amplification factor. Percentage scintillation is the parameter which is normally used for inferring the r_0 values. In order to carry out the calibration of percentage scintillation we are in the process of obtaining simultaneous measurements for r_0 using high resolution H_α observations (Goode et al. 2000; Sridharan et al. 2004). r_0 values can be obtained by *spectral ratio* technique from a series of short exposure H_α observations. We did an initial comparison of our results with already existing calibration plots found in the literature (Beckers 2001). According to these plots corresponding values of r_0 plotted in Fig. 3 ranges from 2.5 to 4.4 for percentage scintillation of 0.165 to 0.095, respectively.

Our future plans include continuous scintillation measurements of MAST site, cross calibration of percentage scintillation with our own r_0 measurements, and deployment of identical calibrated scintillometer in other sites.

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Spatial Damping of Linear Compressional Magnetoacoustic Waves in Quiescent Prominences

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Abstract. We study the spatial damping of magnetoacoustic waves in an unbounded quiescent prominence invoking the technique of MHD seismology. We consider Newtonian radiation in the energy equation and derive a fourth order general dispersion relation in terms of wavenumber k . Numerical solution of dispersion relation suggests that slow mode is more affected by radiation. The high frequency waves have been found to be highly damped. The uncertainty in the radiative relaxation time, however, does not allow us to conclude if the radiation is a dominant damping mechanism in quiescent prominence.

Key words. Sun: prominences—oscillations—magnetohydrodynamics radiation: thermal.

1. Introduction

Prominences are masses of relatively cool ($T \sim 10^4$ K) and dense material suspended in the corona ($T \sim 10^6$ K). The spectra of prominences hold the key to understanding the physical conditions, e.g., temperatures, densities, pressures, etc. The internal structure and physical properties of prominences, however, can be studied through a new tool of prominence seismology. Magnetohydrodynamic (MHD) waves and oscillations of the solar prominence have been carried out both from the ground and from space (Patsourakos & Vial 2002). Small amplitude waves (or oscillations) with velocity amplitudes from 0.1 km s^{-1} to $2\text{--}3 \text{ km s}^{-1}$ have been observed (e.g., Bashkirtsev & Mashnich 1984; Molowny-Horas *et al.* 1999). The velocity field oscillations are classified into three main categories: short-period oscillations ($8 \leq t \leq 20$ min), intermediate-period class ($10 \leq t \leq 40$ min), and long period oscillations ($40 \leq t \leq 90$ min). A very short period oscillation of 30 s has been observed simultaneously from two telescopes (Balthasar *et al.* 1993).

Using the VTT telescope at Sac Peak, Molowny-Horas *et al.* (1999) found velocity perturbations with periods between 28 and 95 min at different locations in a prominence and observed that the amplitude of the oscillations decreases in time with damping times between 101 and 377 min. Terradas *et al.* (2002) investigated the temporal and spatial variations of oscillations and reported the strong damping of oscillations with damping times between two and three times the wave period. Terradas *et al.* (2001) considered the Newtonian radiation in the energy models and found only slow

mode waves being affected by damping leaving fast mode waves almost undamped. Carbonell *et al.* (2004) studied the damping of MHD waves in an unbounded medium by considering the effect of prominence-corona transition region (PCTR) and neglecting the adiabaticity. Terradas *et al.* (2005) investigated the damping of MHD waves in an inhomogeneous and bounded medium by neglecting the adiabaticity. They found out that only slow mode waves are affected leaving fast mode waves almost undamped. This is in agreement with Terradas *et al.* (2001). Recently, Ballai (2003) has reviewed some of the possible mechanisms that can work in the prominence to explain the spatial damping of linear compressional waves.

In this paper, we consider the Newtonian radiation to discuss the spatial damping of MHD waves. MHD equations and dispersion relation are presented in section 2; results and discussions are given in the last section.

2. MHD equations and dispersion relation

Considering a homogeneous equilibrium configuration, which is unbounded in all directions with magnetic field in the x -direction and neglecting the effect of gravity, we have

$$p_0 = \text{const } t, \quad \rho_0 = \text{const } t, \quad T_0 = \text{const } t, \quad \mathbf{B}_0 = B_0 \hat{x}, \quad \mathbf{v}_0 = 0, \quad (1)$$

where p_0 , ρ_0 , T_0 and B_0 are equilibrium values of pressure, density, temperature and magnetic field.

The relevant MHD equations are given by:

$$\frac{D\rho}{Dt} + \rho \nabla \cdot \mathbf{v} = 0, \quad (2)$$

$$\rho \frac{D\mathbf{v}}{Dt} = -\nabla p + \frac{1}{\mu} (\nabla \times \mathbf{B}) \times \mathbf{B} + \rho \mathbf{g}, \quad (3)$$

$$\frac{\rho^{\hat{\gamma}}}{\hat{\gamma} - 1} \frac{D}{Dt} \left(\frac{P}{\rho^{\hat{\gamma}}} \right) = 0, \quad (4)$$

$$\frac{\partial \mathbf{B}}{\partial t} = \nabla \times (\mathbf{v} \times \mathbf{B}), \quad (5)$$

$$\nabla \cdot \mathbf{B} = 0, \quad (6)$$

$$p = \rho RT, \quad (7)$$

where $\hat{\gamma}$ is the complex ratio of specific heats.

We take perturbation in the MHD equations as $\exp i(\omega t + k \cdot r)$ and ω is the frequency of oscillations. The effect of Newtonian radiation can be incorporated by keeping the form of the energy equation adiabatic with complex $\hat{\gamma}$ (Bunte and Bogdan 1994), which is given by

$$\hat{\gamma} = \frac{1 + i\omega\tau_R\gamma}{1 + i\omega\tau_R}. \quad (8)$$

We can generalize the ideal MHD case by replacing sound speed C_s by complex $\hat{C}_s$, which is given by

$$\hat{C}_s^2 = \frac{\hat{\gamma} p_0}{\rho_0}, \quad (9)$$

where $\hat{\gamma} = \gamma$ in the adiabatic limit and $\hat{\gamma} = 1$ in the isothermal limit.

Linearising the MHD equations (2–7), we obtain a fourth order dispersion relation as

$$a_4 k^4 + a_3 k^3 + a_2 k^2 + a_1 k + a_0 = 0, \quad (10)$$

where the coefficients in the dimensionless form are given by

$$\begin{aligned} a_0 &= \omega^5 - \frac{i\omega^4}{\tau_R}, \\ a_1 &= 0, \\ a_2 &= -\omega^3(1 + V_A^2) + \frac{i\omega^2}{\tau_R} \left(\frac{1}{\gamma} + V_A^2 \right), \\ a_3 &= 0, \\ a_4 &= \cos^2 \theta V_A^2 \omega - \frac{i \cos^2 \theta V_A^2}{\gamma \tau_R}. \end{aligned}$$

For spatial damping, we take ω to be real and k to be complex as $k_R + ik_I$. The dispersion relation has four roots, which reveal the nature of magnetoacoustic modes due to damping. The roots of the dispersion relation have been obtained numerically by making use of the physical parameters in prominences (Terradas *et al.* 2001), namely $B_{0x} = 5G$, $\rho_0 = 1.2 \times 10^{-10} \text{ kg m}^{-3}$, $T = 7 \times 10^3 \text{ K}$. We have calculated damping per wavelength, $D_L (= k_I/k_R)$ for different cases.

3. Results and discussion

The real and imaginary parts of complex wavenumber give damping per wavelength, D_L for slow and fast modes. Slow mode is not present when there is no horizontal wavenumber. The dispersion relation gives higher damping per wavelength for slow mode wave compared to fast mode. For a given value of frequency ω , the damping per wavelength first increases as a function of τ_R and attains some maximum value then decreases as we vary τ_R from 10^{-5} to 10^{+5} (in dimensionless units) (Fig. 1). As the radiative relaxation time increases, the damping per wavelength increases which shows greater wave damping at low values of radiative time and then attains some maximum value and then decreases.

The slow mode wave has higher values of damping per wavelength, showing higher levels of damping due to radiation. For $\tau_R \rightarrow \infty$, the wave takes infinite time to damp and therefore travels very long distances.

The value of radiative relaxation time in prominences is not known. According to Terradas *et al.* (2001), the radiative relaxation time could be between 10^{-4} and 10^4 .

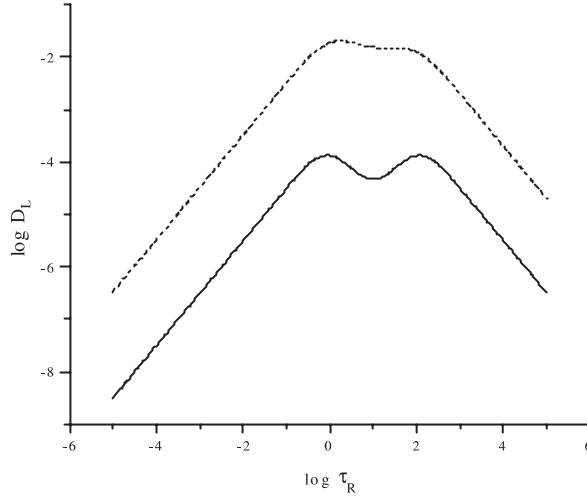


Figure 1. Damping per wavelength, $D_L (= \text{Im}(k)/\text{Re}(k))$ for the magnetoacoustic modes with $\omega = 10^{-1}$ (dotted line: slow mode; solid line: fast mode).

The radiative relaxation time in PCTR varies from 10 – 10^3 s (Terradas *et al.* 2001). Taking the radiative term as

$$L_r = \rho^2 \chi^* T^\alpha,$$

the radiative relaxation time τ_R is given by

$$\tau_R = \frac{\gamma p}{(\gamma - 1) \chi^* \rho^2 T^\alpha},$$

which for different prominence regimes mentioned in Carbonell *et al.* (2004) is between 10^2 and 10^4 s both for prominence and PCTR. The theoretical value of radiative relaxation time in the corona typically is 6×10^4 seconds (Priest *et al.* 1991). The value of radiative relaxation time calculated using opacity values of $(4 \times 10^3 \text{ cm}^2 \text{ g}^{-1})$. Weigert & Wendker (1996) are 1.1×10^{-4} s (Ballai 2003). For a given value of radiative relaxation time, the damping per wavelength increases almost linearly with frequency (Fig. 2). This shows that damping of magnetoacoustic waves could be directly proportional to frequency and higher frequency waves could be damped more in quiescent prominences.

The magnetoacoustic waves can be spatially damped by a few more dissipative mechanisms such as classical viscosity/resistivity, magnetic diffusivity, and thermal conduction and electrical conduction. The kinematic viscosity/resistivity gives very high damping lengths, which shows that the effect of classical viscosity in the damping of magnetoacoustic waves is very small. Using ion collisional rate and Larmor frequency, Ballai (2003) concluded that the nature of viscosity could be isotropic and that of thermal conduction could be anisotropic. Thermal conduction can affect the slow mode waves provided the wavelengths are short.

Singh (2006), studied the spatial damping of the magnetoacoustic waves using the energy losses through Newtonian cooling and MHD turbulence. It was found that the slow mode wave has shorter damping length compared to the fast mode wave. From

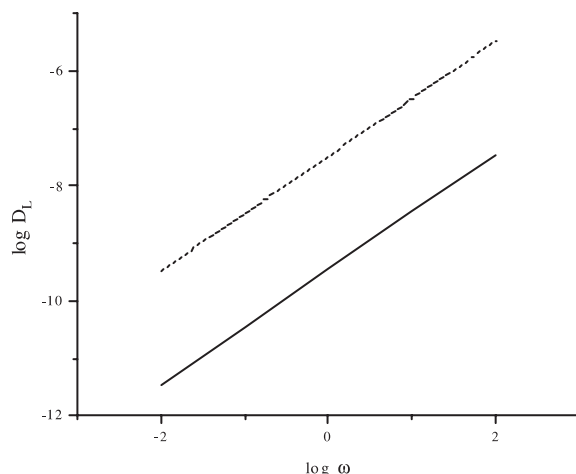


Figure 2. Damping per wavelength, $D_L (= \text{Im}(k)/\text{Re}(k))$ for the magnetoacoustic modes with $\tau_R = 10^{-7}$ (dotted line: slow mode; solid line: fast mode).

prominence seismology, the opacity and turbulent kinematic viscosity is inferred. It is found that the turbulent viscosity is orders of magnitude higher than the Spitzer's one. Also, it has been found that the calculated turbulent viscosity can reproduce the observed damping time and damping length in prominences. The convective disturbances may travel up in the solar atmosphere through the magnetic field lines that are anchored in the photosphere and cause the MHD turbulence in prominences.

In conclusion, I find that the Newtonian radiation alone is inadequate to explain the spatial damping of both slow and fast mode waves and gives acceptable damping lengths for certain values of radiative relaxation time. The presence of density inhomogeneity may lead to phase mixing and resonant absorption of MHD waves. The uncertainty in the radiative relaxation time in prominences does not allow us to reach a definite conclusion as to whether the oscillations are damped by Newtonian radiation alone. In order to explain the damping of both slow and fast modes, some additional mechanisms such as turbulent viscosity, ion-neutral damping, wave leakage, phase-mixing, resonant absorption, etc. are worth considering.

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Observational Searches for Chromospheric g-Mode Oscillations from CaII H-Line Observations

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Abstract. We have used a high spatial and temporal resolution of long time sequence of spectra in CaII H-line obtained at the Vacuum Tower Telescope (VTT) of the Sacramento Peak Observatory on a quiet region at the center of the solar disk over a large number of bright points and network elements to search for atmospheric (chromospheric) g-mode oscillations. An important parameter of the H-line profile, intensity at H_{2V} ($I_{H_{2V}}$), has been derived from a large number of line profiles. We derived the light curves of all the bright points and network elements. The light curves represent the main pulse with large intensity amplitude and followed by several follower pulses with lower intensity amplitudes. The light curves of these bright points would give an impression that one can as well draw curves towards and away from the highest peak (main pulse) showing an exponential growth and decay of the amplitudes. An exponential decaying function has been fitted for all the light curves of the bright points to determine the damping time of the modes that are more or less the same, and one value of the coefficient of exponent can represent reasonably well the decay for all the cases. The FFT analysis of temporal variation of both the bright points and the network elements indicates around 10-min periodicity. We speculate that this longer period of oscillation may be related to chromospheric g-mode oscillations.

Key words. Solar chromosphere: CaII H-line—oscillations chromospheric g-mode.

1. Introduction

It is known that there are two types of gravity waves in the Sun, namely, (i) the internal gravity waves, that are expected to be confined to solar interior, and (ii) the atmospheric gravity waves, which are related to the photosphere, chromosphere, and may be further beyond. Pallé (1991) had discussed in great detail the various methods to search for solar gravity modes. Many authors have claimed, in the past 22 years, to detect internal g-modes in the Sun, but so far, there is no observational evidence. The atmospheric g-modes of small scale may exist, but in the atmosphere the Brunt Vaisala frequency is ≈ 5 mHz and hence one would expect the observed frequencies to be a few mHz. And of course, there would be some very high order atmospheric

g-modes with longer periods. There are some observational evidences to show that there is signature of atmospheric gravity waves at the chromospheric level using the time sequence of filtergrams obtained in CaII K and Mg b2 lines (Damé *et al.* 1984; Kneer & von Uexkull 1993).

In this paper, we make an attempt to search for atmospheric (chromospheric) g-modes using the intensity oscillations at the sites of the chromospheric bright points and network elements observed under a high spatial, spectral and temporal resolution in CaII H-line.

2. Data and analysis

The observations were obtained by Sivaraman with the Vacuum Tower Telescope (VTT) using the echelle spectrograph of the Sacramento Peak Observatory on September 13, 1971 under the Program B of the HIRKHAD mode (Beckers *et al.* 1972). We have totally 177 frames for the 35-min duration of observations with a repetition rate of 12 s. We have chosen 29 bright points and 3 network elements and designated them as $B_1, B_2, B_3, \dots, B_{32}$ (refer Fig. 1 of Kariyappa *et al.* 1994, hereafter KSA). We have derived photometrically calibrated line profiles for all the bright points and network elements. We have measured an important parameter of the line profile, namely, the intensity of the emission peak on the violet side (I_{H2V}) for each profile, and plotted I_{H2V} versus time. We have done the power spectrum analysis using FFT for the time series data of the bright point (B_1) and the network (B_4).

3. Results and discussion

In our earlier paper (KSA), we have discussed on the classification of bright points and shown that the bright points can be grouped into three classes (see KSA, Figs. 2 and 3: class I; Fig. 4: class II and Fig. 5: class III) depending on their intensity enhancements

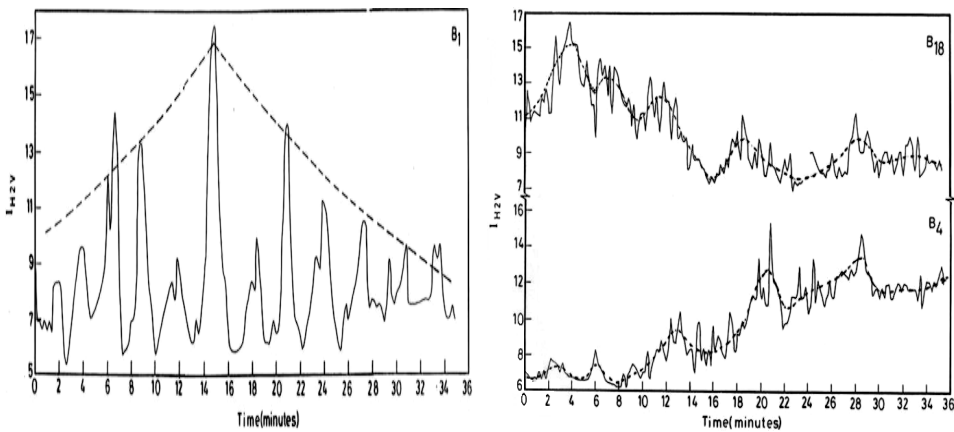


Figure 1. Left box: The light curve of a bright point B_1 . The thick and dotted lines represent the 3-min oscillations and an exponential decaying in brightness respectively. Right box: The light curves of the two network elements B_4 and B_{18} . The dotted line represents the mean curve of the intensity fluctuations showing 5–7 min periodicity.

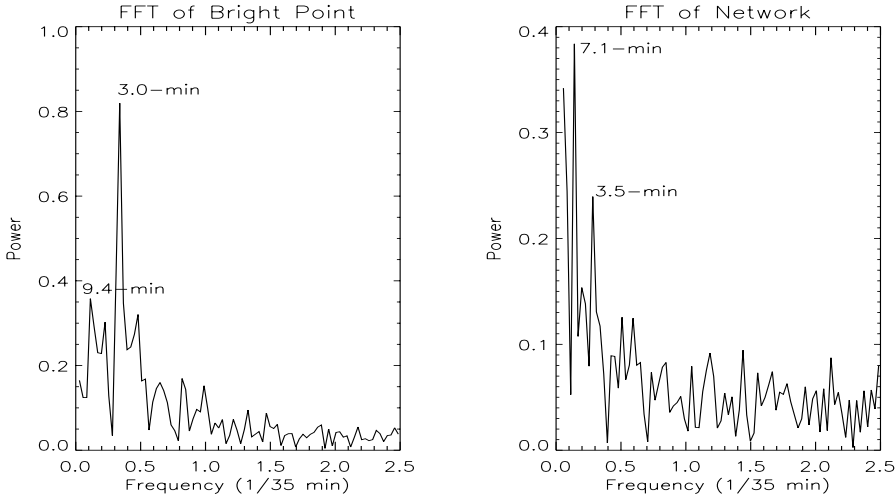


Figure 2. The power spectra of bright point (B_1 , left box) and of network (B_4 , right box).

during their dynamical evolution. A majority of the class I bright points show a large enhancement of $I_{H_{2V}}$ at their peak brightness phase, as high as 3 times above the mean ambient level (refer Figs. 2 and 3 of KSA), which corresponds to the undisturbed line profile (at time $t = 0$) of Fig. 11 of KSA. Class II bright points show moderate intensity enhancement in $I_{H_{2V}}$ (about twice the mean ambient level) at the peak brightness phase. However, class III bright points show only a marginal increase in $I_{H_{2V}}$ at the brightest phase. All these results suggest that the different classes of bright points may be associated with different strengths of the magnetic field. In the light curves of the bright points, the highest peak in the intensity variation of each bright point is designated as the “main pulse”, and marked as P_1 (refer Figs. 2 to 5 of KSA). The main pulse is followed by several pulses with smaller amplitudes and this would give an impression that one can as well draw curves towards the highest peak showing an exponential growth of the amplitudes. We suggest the possibility that the main pulse is followed by several pulses whose amplitudes decay exponentially with time.

In order to know the damping time of the modes, we have fitted exponential function for all the classes of bright points using time series data. The slopes of decay of the curves are more or less the same, and mean value of coefficient of the exponent ($4.14 \times 10^{-4} \text{ s}^{-1}$) taken for all the 29 bright points can represent reasonably well the decay for all the cases. We have examined the light curves of the bright points obtained by Liu (1974, see Fig. 6) and find that there is a strong main pulse and it is followed by several pulses with exponentially decreasing in their intensity amplitudes. Similarly we have fitted the exponential functions for his light curves of bright points (see his Fig. 6) and found that the slopes of decay are the same for any bright points and the value of coefficient of exponent is found to be $3.5 \times 10^{-4} \text{ s}^{-1}$. It is clearly seen that there is a good agreement between the coefficient of exponent values derived from both Liu’s and the present data. The value of coefficient of exponent represents the damping time of the modes and it is of the order of 40 minutes. For comparison, we have tabulated the exponential functions for the light curves of B_1 and B_5 bright points and for Liu’s (1974) bright points in Table 1.

Table 1. Exponential decay in the amplitudes of the follower pulses from the main pulse for the bright points (B_1 and B_5) and Liu's data.

Bright point	Exponential function; $I(t)$
B_1	$16.69 \exp(-5.9 \times 10^{-4}t)$
B_5	$18.89 \exp(-4.9 \times 10^{-4}t)$
Liu bright point 1	$13.95 \exp(-3.4 \times 10^{-4}t)$
Liu bright point 2	$15.47 \exp(-3.6 \times 10^{-4}t)$

We have shown the light curve for the bright point B_1 (not shown for B_5) in the left side box of Fig. 1. It is clearly seen from Fig. 1 (left box) that the bright point shows a 3-min period of intensity oscillations. In the left side box of Fig. 1, the thick line represents the 3-min period of intensity oscillations, whereas the dotted line shows an exponential decaying in brightness. In addition to the light curves of the bright points, we have generated the light curves for the two network elements (B_4 , B_{18}) from the same time sequence spectra, and these are shown in the right side box of Fig. 1. We know from the earlier studies (KSA, Kariyappa 1994; Kariyappa *et al.* 2005) that the network elements have different patterns of intensity oscillations showing 5–7 min periodicity (dotted curve in Fig. 1, right box). We have done the power spectrum analysis on the time series data of both bright point and network to look for longer periods of oscillations. In Fig. 2, we have presented the power spectra for B_1 in the left panel and for B_4 in the right panel. One can see from the power spectrum that there is an indication of a longer period of oscillation around 10-min.

We can summarize the preliminary results as follows:

- The chromospheric bright points and network elements exhibit different periodicities (3-min and 5–7 min respectively), but, both of them seem to be also associated with the 10-min longer period of oscillations.
- The light curves of the bright points show an exponential growth and decay of the amplitudes and the damping time of the modes is of the order of 40-min.
- We speculate that the 10-min longer period of oscillations may be related to g-mode oscillations of the chromosphere. These results confirm the earlier findings that there is a signature of gravity waves in the chromosphere (Damé *et al.* 1984; Kneer & von Uexkull 1993) that needs further investigation with a long stretch of observations.

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The Extreme Solar Activity during October–November 2003

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Abstract. The positional measurements of sunspots from the Kodaikanal Observatory and Solar Geophysical data are used to study the association between occurrence of the abnormal activities of big sunspot groups that were observed during the period of October–November 2003 and occurrence of the flares. During the evolution of the sunspot groups, we have investigated the temporal variations in (i) areas; (ii) rotation rates; (iii) longitudinal extents; and (iv) number of small spots produced in a sunspot group. Among all these activity variations, we find that the spot groups that experience abnormal rotation rates during their evolutionary phases eventually trigger the flares.

Key words. Sunspots—rotation—flare—extreme activity.

1. Introduction

The Sun influences the Earth's environment significantly (Hiremath & Mandi 2004 and references therein) that requires constant observations and understanding the solar activity phenomena. The solar activity is mainly caused by the sunspots that are supposed to be tracers of subsurface dynamics (Howard 1996; Hiremath 2002 and references therein).

Using Kodaikanal Observatory daily white light pictures and H α flare data, we (Hiremath & Suryanarayana 2003) have shown that either leading or following or both of the bipolar spots that experience abnormal rotation rates eventually trigger the flares. Since we used the daily data we were unable to pinpoint exactly, after sunspot experiencing the abnormal rotation rates, at what time of the day the flare occurs. Moreover, it is interesting to know whether the sunspots with complex magnetic activity (considered in the present analysis) also experience the abnormal rotation rates and eventually produce the flares. In the present analysis, we use high temporal resolution (varies from ~ 30 minutes to ~ 10 hours) of the positional measurements of the sunspot groups that occurred during October–November 2003.

2. Data and analysis

We are interested to study the extreme activities of the sunspot groups (10484, 10486 and 10488 according to NOAA classification) that occurred during the period of

October–November 2003 and produced many flares and CMEs. The positional measurements of the sunspot groups and the occurrence of the H α flares are considered from the archive of Solar Geophysical data (<http://sgd.ngdc.noaa.gov/sgd/jsp/solarindex.jsp>) that is contributed from different observatories all over the world. In order to have better statistics, we add the data of the sunspot positional measurements obtained from the Kodaikanal Observatory to the data of Solar Geophysical sunspot positional measurements. The data of positional measurements consists of the time of observations, the heliographic latitude, the longitude from the central meridian, the area (A), the number of small sunspots in a group (NS) and the longitudinal extent (LE).

We compute rotation rate ω_i of the sunspot groups as follows:

$$\omega_i = \frac{(l_{i+1} - l_i)}{(t_{i+1} - t_i)}, \quad (1)$$

where l is the heliographic longitude from the central meridian, t_i is the time of observation, $i = 1, 2, 3, \dots, n$, and n is the *lifespan* of the spot group.

The typical white light images on 29 and 30 October 2003 that are taken from the Kodaikanal Observatory are presented in Fig. 1. The different activity parameters such as A (millionth of hemisphere), NS , LE (degrees) and ω (deg/day) are presented in Fig. 2. In Fig. 2, the blue continuous line is variation of different activity parameters for the sunspot group number 10484, the red dotted line for the group number 10486 and, the green dashed line for the group number 10488 respectively. In the same figure the red continuous vertical bars represent the occurrence dates of the flares. The dates along the x axis are counted from the beginning of January 1. For example, on January 1, the date is counted as 1; February 1 as 32 and so on. Thus for the sunspot activity 10484 that occurs on October 18, the date along the x axis starts from 290.

We define *abnormal activity* as follows: For each spot group and for all the time of observations, first we compute the mean ($\bar{x}$) and the standard deviation (σ) of the respective activity parameters: A_i , NS_i , LE_i and ω_i ; where $i = 1, n$ and n representing the number of time of observations (for example $i = 1, 37$ for the spot group 10484, $i = 1, 39$ for the spot group 10486 and $i = 1, 29$ for the spot group 10488 respectively) during the lifespan of a spot group. For each time of observations, if absolute value of the difference $(\bar{x} - x_i) > 1\sigma$ (where $\bar{x}$ is the mean and x_i represent respective parameters as defined above), then we consider the corresponding parameter at that time as *abnormal activity* of the spot group. For example, for the spot group 10484 that has 37 number of observations during its 12 days lifespan, $\bar{A}$ is the mean area and A_i are the area values of the sunspot group for i number of observations. For each A_i value, we compute the difference $(\bar{A} - A_i)$ and if its absolute value is $> 1\sigma$, then we consider the abnormal activity of area of the sunspot group on the particular time of observation.

3. Results and discussion

A typical data set that relates the occurrence numbers of the abnormal activity of different parameters and correspondingly the occurrence numbers of the flares for each day is presented in Figs. 3 and 4 respectively.

In order to know the association between the occurrence of the abnormal activity and occurrence of the flares, we compute the correlation coefficient and its significance.

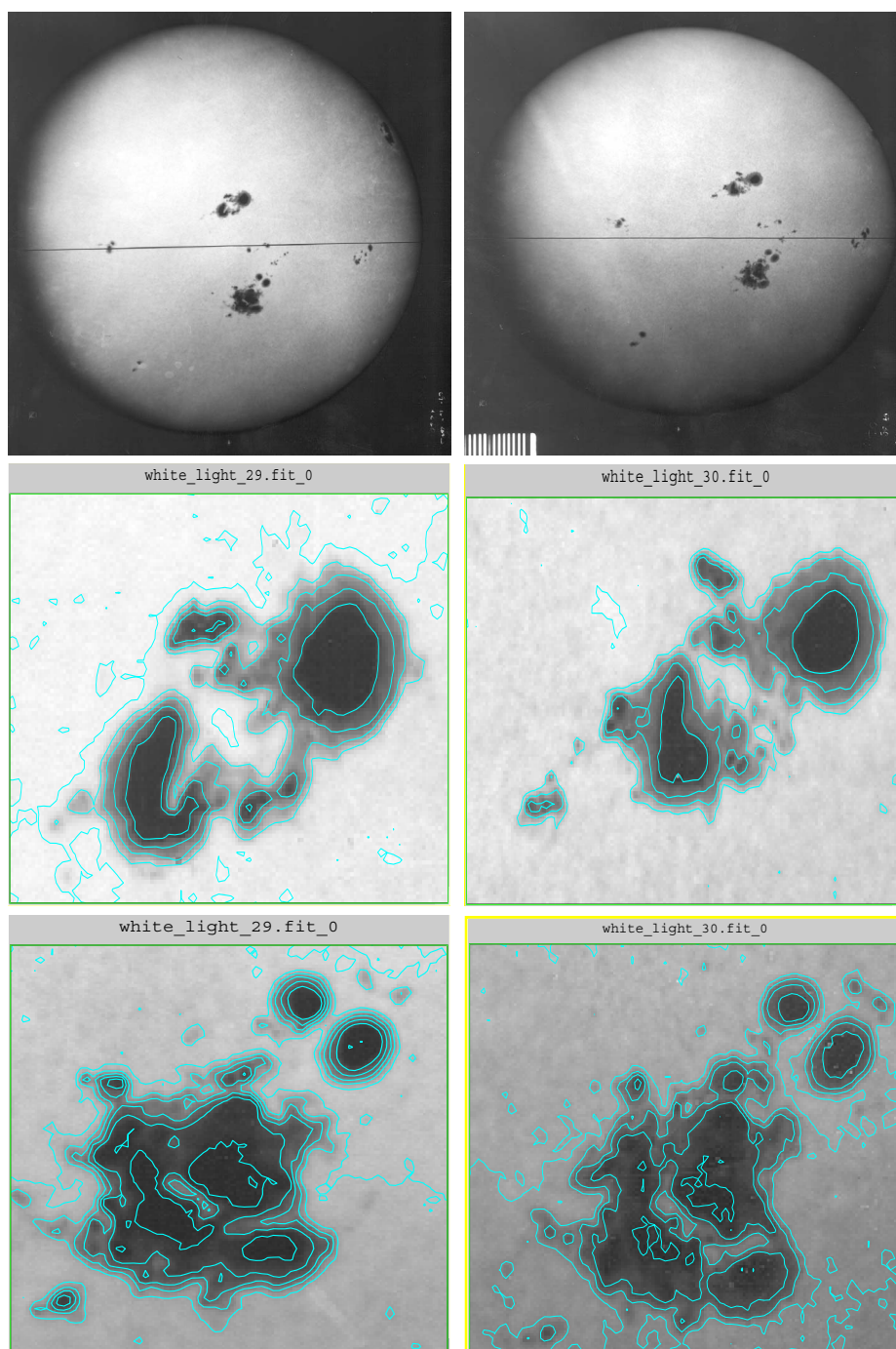


Figure 1. The Kodaikanal white light images observed on 29 and 30 October 2003 respectively. **Top panel:** Full disk images. **Middle panel:** The big sunspot group 10488 in the northern part of the solar disk that is close to the equator. **Lower panel:** The big sunspot group 10486 in the southern part of the solar disk that is close to the equator. Different levels of intensity (decreasing from penumbra region to umbra region) contours are superposed on the sunspots.

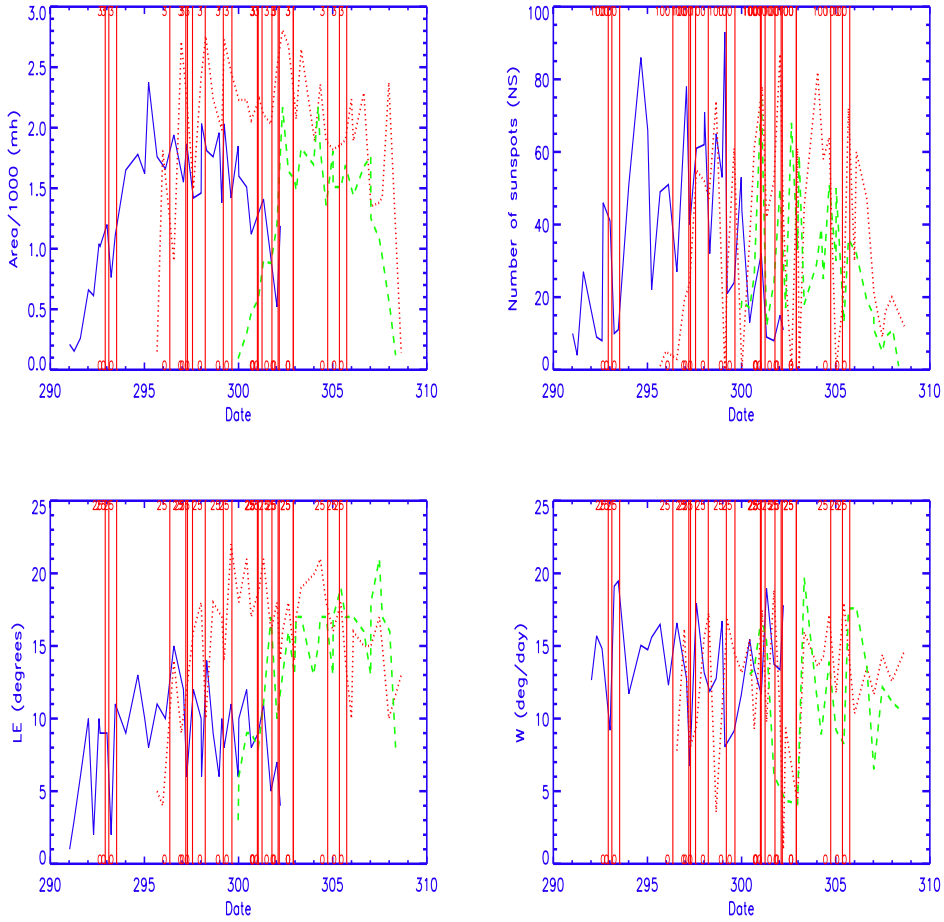


Figure 2. The variation of different activity parameters such as area (in millionths of hemisphere), number of sunspots, longitudinal extent (in degrees) and rotation rates (in deg/day) of big sunspot groups that occurred during October 18 to November 4, 2003. The lines in different colours are as follows: the blue continuous line represents the sunspot group number 10484, the red dotted line represents the number 10486 and the green dashed line represents number 10488. The red vertical bars represent the occurrence of dates of the flares.

When we say the correlation coefficient, we mean the Spearman Rank Correlation coefficient (Press *et al.* 1992). This method of finding the correlation between two variabilities is more robust than the correlation found from the usual method. We know (Fisher 1930; Elmore & Woehlke 1996), while deriving the correlation coefficient, that both the variabilities must be linearly related and we should get a good least square fit for the linear relationship of the form $F = A + BS$, where S is occurrence number of abnormal activity and F is occurrence number of the flares, A and B are coefficients in the linear equation. If the correlation is very good, points of both the variabilities should fall closer to the straight line (in red colour for all the figures presented in our study) obtained from the least square fit and significance of the correlation coefficient should be very high.

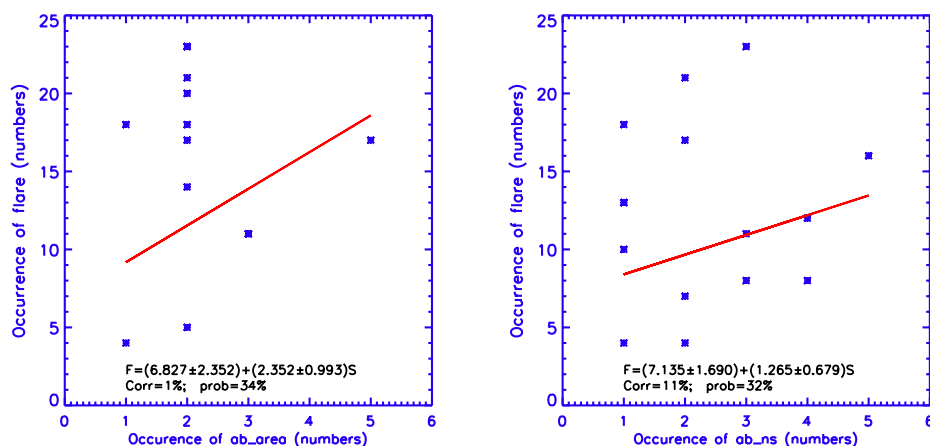


Figure 3. The abnormal activity of the sunspot groups *versus* occurrence of the flares. The two figures represent occurrence of the abnormal area and occurrence of the abnormal number of sunspots in a spot group with respect to the occurrence of the flares. The red continuous straight line is obtained from the linear least square fit of both sets of the data. Here S represents occurrence of the respective abnormal activity of either area or the number of sunspots and F is the occurrence of the flares.

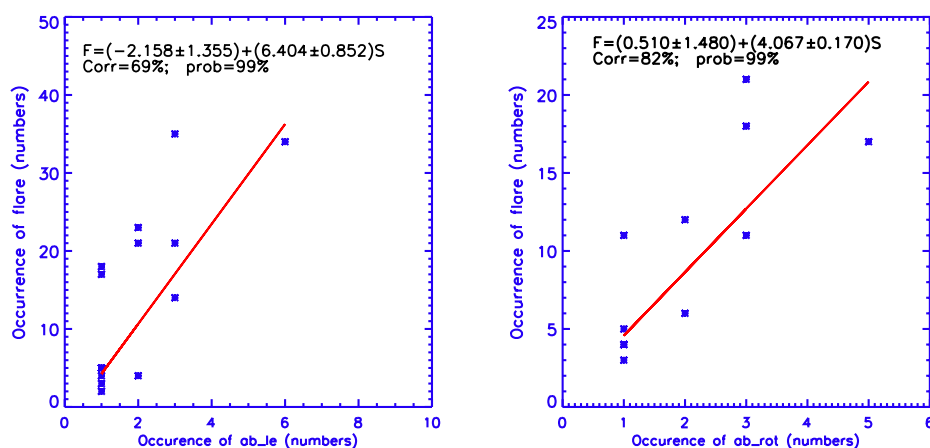


Figure 4. The abnormal activity of the sunspot groups *versus* occurrence of the flares. The two illustrations represent the occurrence of the abnormal variations in the longitudinal extent and the rotation rates of the sunspot groups with respect to the occurrence of the flares. The red continuous straight line is obtained from the linear least square fit of both sets of the data. Here S represents occurrence of the respective abnormal activity of either longitudinal extent or rotation rates and F is the occurrence of the flares.

It is interesting to note from Fig. 4 (right) that, compared to other abnormal activities, the occurrence of abnormal rotation rates of the sunspot groups show the strong association (correlation coefficient is $\sim 82\%$ with a very high probability) with the occurrence of the flares. This implies that the unknown perturbations along the longitudinal direction below the surface is one of the main causes for triggering the flares. This

view can also be corroborated from the strong association between the occurrence of abnormal activity of the variation of the longitudinal extent of each spot group and the occurrence of the flares.

Presently, it is believed that the flares occur due to magnetic reconnection wherein oppositely directed magnetic fluxes merge and annihilate each other producing the required amount of flare energy. We don't have any conclusive evidences whether reconnection takes place either below or above the photosphere. There are some studies (Zirin & Liggett 1987; Leka et al. 1996; Canfield & Pevtsov 2000; Hiremath & Suryanarayana 2003; Hiremath et al. 2005; Hiremath 2005) suggesting that reconnection events may be taking place below the photosphere. In both the cases of reconnection events that may be occurring above or below the surface, approaching of sunspots towards each other and then merging of foot points is necessary. From the observed data, we find that during the course of their evolution the three spot groups initially appear to be bipolar and then evolve as complex beta-gamma-delta magnetic regions. It is interesting to note that except the spot group 10488, most of the very high energy ($> X10$) flares occur during late stages of their life times. This suggests that the flux tubes while raising towards the surface, experience inequilibrium at certain regions of the convective envelope eventually bringing together the flux tubes of different polarities in a spot group that ultimately trigger the flare. Inequilibrium of the spot groups can be attained by either strong flows or rotational gradients below the surface (Hiremath & Suryanarayana 2003). In fact the helioseismic inferences show both strong flows (Zhao et al. 2001, 2004) and rotational gradients below the photosphere (around 0.935 radius of the Sun). From this study, we cannot estimate the depth of initial anchoring of these spot groups. Thus both the possibilities are not ruled out for the loss of the equilibrium of the clustered flux tubes in a spot group that triggered the flares during October–November 2003.

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Study of Microflares through SOXS Mission

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Abstract. We present a study of 10 microflares observed in 4–30 keV by SOXS mission simultaneously with H α observations made at NAOJ, Japan during the interval between February and August 2004. The X-ray and H α light curves showed that the lifetime of microflares varies between 4 and 25 min. We found that the X-ray emission in all microflares under study in the dynamic energy range of 4–30 keV can be fitted by thermal plus non-thermal components. The thermal spectrum appeared to start from almost 4 keV, low level discriminator (LLD) of both Si and CZT detectors, however it ends below 8 keV. We also observed the Fe line complex features at 6.7 keV in some microflares and attempted to fit this line by isothermal temperature assumption. The temperature of isothermal plasma of microflares varies in the range between 8.6 and 10.1 MK while emission measure between 0.5 and $2 \times 10^{49} \text{ cm}^{-3}$. Non-thermal (NT) emission appeared in the energy range 7–15 keV with exponent $-6.8 \leq \gamma \leq -4.8$. Our study of microflares that had occurred on 25 February 2004 showed that sometimes a given active region produces recurrent microflare activity of a similar nature. We concluded from X-ray and simultaneous H α observations that the microflares are perhaps the result of the interaction of low lying loops. It appears that the electrons that accelerated during reconnection heat the ambient coronal plasma as well as interact with material while moving down along the loops and thereby produce H α bright kernels.

Key words. X-ray detectors—coronal 100 ps—solar flares—microflares—hard X-ray emission.

1. Introduction

Hard X-rays ($> 20 \text{ keV}$) were discovered with a balloon-borne instrumentation to occur on an average once every $\sim 6 \text{ min}$ near solar maximum (Lin *et al.* 1984). This investigation considered that the energy released in accelerated electrons, summed over HXR bursts of all sizes, might contribute significantly to the heating of the active corona. Recently, Lin *et al.* (2001) investigated HXR microflares up to 8 keV using BATSE/CGRO detectors. They found that the generally steep HXR spectra (power-law fits with exponent 3–7) reveal that most of the flare energy is in the non-thermal

electrons at lowest energy. Furthermore, the similarity between the distributions of frequency vs. energy release for > 8 keV bursts and the active region transient brightening (ARTBs) seen in soft X-rays by the SXT/YOHKOH instrument (Shimizu 1995) suggested that these accelerated electrons may provide the energy for ARTBs. Recently Liu *et al.* (2004) have reported that microflares are seen in $H\alpha$, soft and hard X-ray wavelengths and their temporal evolution resembles large flares. They found that these small impulsive microflare events are located near well-defined magnetic neutral lines in active regions. The X-ray spectral analysis of the microflares and their association with type III radio bursts could enable them to conclude that the X-rays over 10 keV are emitted by non-thermal electrons.

Sylwester *et al.* (2000) used gas filled proportional detectors to improve sensitivity and spectral resolution in the energy band 2 keV to 8 keV, which, however, revealed FWHM > 1 keV at 5.9 keV, but succeeded in observing microflares. RHESSI (Lin *et al.* 2002) uses germanium detectors to achieve this goal by cooling them to < 75 K in order to reduce the dark counts (noise) significantly. On the contrary, our SOXS payload (Jain *et al.* 2005) employs near-room temperature Si and CZT diodes, which operate between -5 and -30°C with the help of thermoelectric cooler coupled with the detector itself. These detectors (Jain *et al.* 2000, 2002a, b) provide high sensitivity for handling a large number of counts and high spectral resolution (sub-keV at < 10 keV). These features of the detectors however, enabled us to observe the X-ray emission from microflares to major flares as a full disk integrated contribution. Thus, in order to study the contribution from individual microflares we have to observe them during the passage of a single active region on the solar disk. Jain *et al.* (2005) described in detail the potential use of SOXS mission to study the microflares.

In this paper, we study microflares using temporal and spectra data from the Si and CZT detectors of the SOXS mission. We aim to address the question of whether the non-thermal nature of microflares extends below 10 keV, currently known as low energy cut-off limit for hard X-ray emission, and whether thermal emission also contributes to the microflares.

2. Observations

Flares on GOES intensity class $\leq C1.0$ are defined as microflares (Chang *et al.* 2004). We also adopt this definition to identify the microflares. Kumar (2004) studied the statistical relationship between SOXS and GOES flares as both missions are in geostationary orbit and both measure full disk integrated X-ray emission. He found correlation coefficient 0.9 between flares observed by these two missions. The best fit of correlation was presented by Kumar (2004) in the form of the following empirical relation:

$$\text{Log [SOXS flux (Y)]} = A + B^* \text{log [GOES flux (X)]}, \quad (1)$$

where SOXS flux is in counts/s and GOES flux is in watts/m^2 . Based on the study of a large sample of flares, Kumar (2004) derived A and B for Si and CZT detectors of SOXS for 7–10 keV energy band as the following:

For Si:

$$A = 8.07,$$

$$B = 1.12,$$

Table 1. List of SOXS microflares under investigation.

Sl. no.	Date	Time(UT)			Peak int.	GOES class	H α class	Location	NOAA region no.
		Begin	Peak	End	7–10 keV CZT				
1	25 Feb 04	04:07:53	04:09:24	04:15:03	17	B5.5	Sf	N13W03	10564
2	25 Feb 04	04:15:03	04:17:06	04:22:04	20	B8.9	Sf	N13W03	10564
3	25 Feb 04	04:24:03	04:25:25	04:29:18	15	B9.2	Sf	N13W03	10564
4	25 Feb 04	05:02:23	05:06:35	05:22:14	30	B9.8	Sf	N14W03	10564
5	25 Feb 04	05:22:14	05:27:06	05:31:58	22	C1.0	Sf	N13W03	10564
6	25 Feb 04	06:45:58	06:51:18	06:56:20	26	C1.0	Sf	N13W03	10564
7	01 May 04	04:33:31	04:43:39	04:58:00	16	B8.9	Sf	S10W25	10601
8	22 Jun 04	05:20:23	05:22:03	05:24:05	34	B8.1	Sf	S13W15	10635
9	09 Aug 04	04:06:39	04:12:43	04:18:58	23	B9.2	Sf	S12E45	10656
10	09 Aug 04	04:18:58	04:25:24	04:31:50	19	B8.2	Sf	S12E45	10656

and for CZT:

$$A = 6.77,$$

$$B = 0.844.$$

The above study enabled Kumar (2004) to define the higher threshold level of microflares for example in CZT detector to set at 60 count/s, i.e., C1 in GOES class. Thus while selecting the microflares for our current study we chose only those flares as microflares, which have been observed by SOXS under the above definition plus either in GOES or in H α . In Table 1, we show the list of microflares observed by SOXS mission and analysed by us together with their details in GOES and H α wavebands. A total of 10 microflares observed by Si and CZT detector of SOXS mission between February and August 2004 were considered for our detailed study. The H α observations were made at National Astronomical Observatory of Japan (NAOJ), Mitaka, simultaneous to the SOXS mission. The seeing was average between 1 and 2 arcsec during all H α observations.

3. Analysis and results

Details of the SOXS mission, in-flight performance, calibration, instrumental response and background are described by Jain *et al.* (2005). The data format and data analysis techniques are also presented by him. SOXS delivers two types of data, viz., temporal and spectral mode. In the temporal mode CZT detector provides light curve in 6–7, 7–10, 10–20, 20–30 and 30–56 keV energy bands. The microflare in any of these energy bands is identified if its signal is $\geq 3\sigma$ than the background in the respective energy band. After unambiguous identification of the microflare it is cross-correlated with GOES and/or H α observations on time and intensity scales. Then we undertake detailed spectra analysis for the period of microflare using standard SOXS software – SOXSsoft (Patel and Jain 2005). In Fig. 1, we show the light curve of the microflares

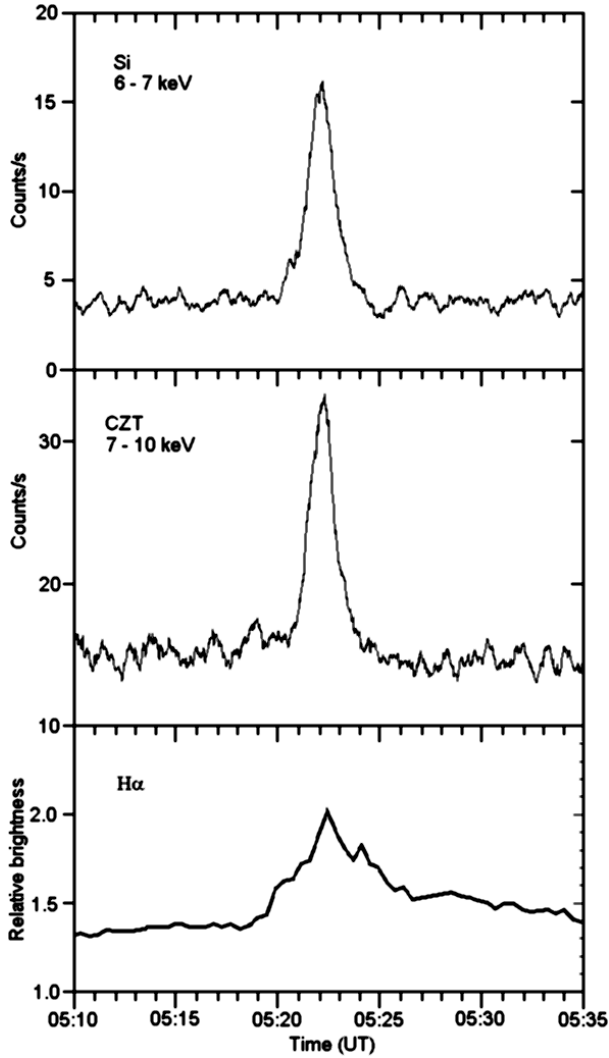


Figure 1. Light curves of 22 June 2004 microflare. **Top and middle panels:** X-ray emission in 6–7 and 7–10 keV as seen by Si and CZT detectors of SOXS mission respectively as a function of time. **Bottom panel:** variation H α relative brightness over the same time interval.

observed on 25 February 2004 in 6–7 and 7–10 keV energy bands by Si and CZT detectors respectively. Also shown in the bottom panel is the light curve of H α relative brightness. In Fig. 2, we show H α filtergrams of the typical microflare observed on 22 June 2004. It may be noted that the two bright kernels are the foot points of the bright loop connecting them. We observed the loop-like structure in the images before the flare, thus apparently the loop-like structure is a part of the pre-existing plage. The brightness of the loop-like structure increased a little bit during the flare, but the increase of the brightness of the kernels was much more remarkable. The six images in Fig. 2 show a remarkable change of brightness of the loop as well as of the kernels. The light curves enabled us to estimate the lifetime of each microflare. We

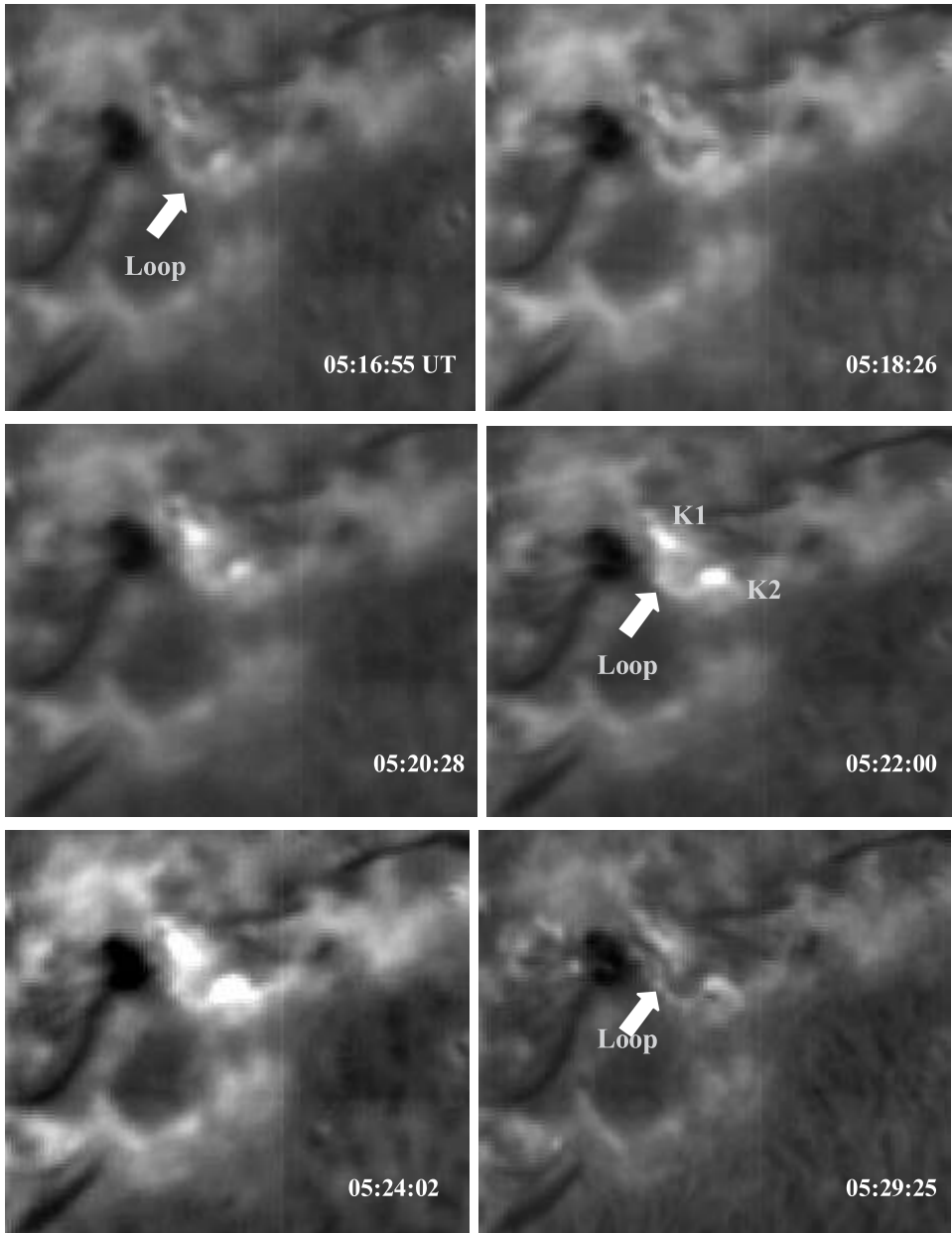


Figure 2. $H\alpha$ filtergrams of 22 June 2004 flare. Two bright kernels K1 and K2 connected by a small bright loop, as shown in the filtergram at 05:22:00 UT, may be noted. The loop and the kernels faded during the decay phase.

found that it varies between 4 and 25 min. In $H\alpha$ these microflares are classified as of Sf importance.

We have carried out a detailed analysis of the X-ray photon spectra of each flare. In Fig. 3, we have shown Si and CZT detectors photon spectra of 22 June 2004

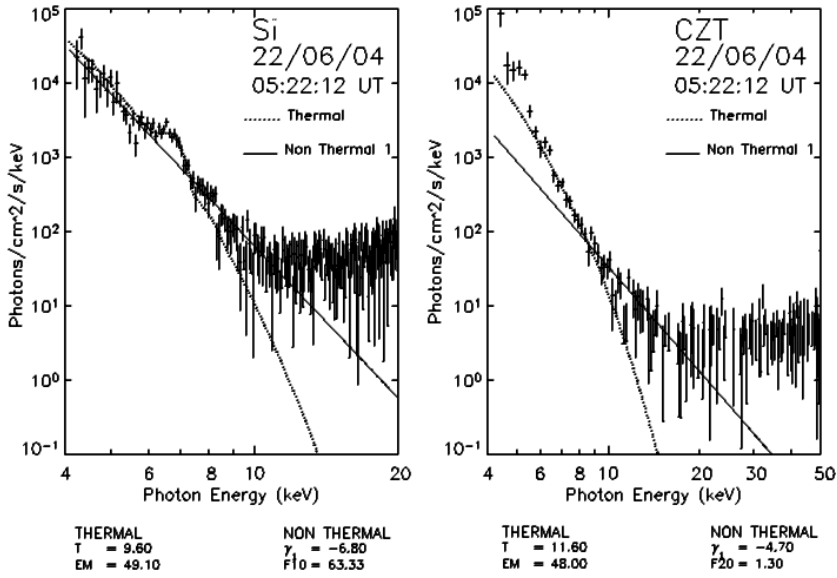


Figure 3. X-ray photon spectra of 22 June 2004 microflare as seen by Si and CZT detectors. The observed photon flux at every channel is represented by $\pm 1\sigma$ error. In Si spectra Fe line complex at 6.7 keV may be noted. CZT shows only continuum X-ray spectrum. Thermal and non-thermal fit is presented for spectra from both detectors. Si detector shows X-ray photon spectrum for the microflare up to 15 keV while CZT up to 25 keV.

microflare that occurred at 05:22:12 UT. The photon spectra are fitted with thermal and non-thermal components. It may be noted from the figure that the photon spectra of this microflare reveal Fe line complex at 6.7 keV. However, Fe/Ni line complex is not visible. On the other hand, our analysis showed that each microflare does not exhibit Fe line complex, which otherwise is observed in all higher important class flares (Jain *et al.* 2005). The isothermal plasma emission was considered to fit the X-ray emission below 10 keV including Fe line complex but we found that the contribution of thermal component was valid below 8 keV in all microflares under study. The thermal spectrum appeared to start from almost 4 keV, i.e., low level discriminator (LLD) of both Si and CZT detectors, however, it ends around 8 keV. The temperature of the isothermal plasma of microflares varies in the range between 8.6 and 10.1 MK while emission measures between 0.5 and $2 \times 10^{49} \text{ cm}^{-3}$ for different microflares studied in this investigation. The non-thermal component was found to begin from 7 keV, and therefore power-law fit between 7 and 14 keV was more appropriate. However, the photon index (exponent) γ varies as $-6.8 \leq \gamma \leq -4.8$ among different microflares.

4. Discussion

A series of microflare activity may be noted from Table 1 in the NOAA active region no. 10564 on 25 February 2004. This active region produced recurrent microflare activity. Almost all microflares produced by this region showed an almost similar kind of X-ray photon spectrum indicating that a single type of energy release mechanism was operating all over the active region. $H\alpha$ filtergrams of the active region

under study revealed that either a thin filament channel or small filament showed considerable activity preceding each microflare. This indicates that microflares are perhaps the result of the interaction of low lying loops. It appears that the electrons accelerated during reconnection caused by the interaction of the low lying loops heat the ambient coronal plasma as well as interact with material while moving down along the loops. The heated coronal plasma produced soft X-ray emission as revealed in the X-ray photon spectrum between 4 and 8 keV including Fe line complex. These accelerated electrons when interacted with chromosphere produced H α bright kernels as shown in Fig. 2. The X-ray photon spectra between 7 and 14 keV showed non-thermal component. We conclude from X-ray photon spectra and simultaneous H α filtergrams that thermal and non-thermal mechanisms are operating in the microflares in agreement with Benz & Grigis (2002); Krucker *et al.* (2002) and Liu *et al.* (2004). The other possibility of microflare occurrence through reconnection of the higher altitude coronal loops and there by accelerated electrons which could heat the plasma at the intersection region of the separatrices of the active region cannot be ignored.

5. Conclusion

We analyzed 10 microflares for the current study and found that the X-ray photon spectra below 8 keV are of thermal nature, while above 8 keV are of non-thermal nature. We conclude that the non-thermal electrons accelerated at the loop apex move along such low lying coronal loops and interact with chromospheric material at the foot points and produce H α bright kernels. The thermal electrons produce soft X-ray emission observed by GOES and SOXS mission in low energy bands.

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Study of a Large Helical Eruptive Prominence Associated with Double CME on 21 April 2001

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Abstract. Here we present a preliminary analysis of a helical eruptive prominence at the east limb of the Sun on 21 April 2001. Unusually this eruption is associated with a double CME. We have tried to study the morphology of the event, energy budget of the prominence and associated CMEs. Our analysis shows that the prominence and first CME started simultaneously from the limb and prominence carries sufficient energy to feed both the CMEs. Moreover, it is also concluded that CMEs are magnetically driven and internally powered.

Key words. Sun: prominence, Coronal Mass Ejection.

1. Introduction

Solar prominences are ribbons of cool (~ 8000 K) dense gas ($\sim 10^{-11}$ gcm $^{-3}$) embedded in the hot tenuous corona, which forms the outer atmosphere of the Sun. Magnetic fields play an important role in supporting the prominence against gravity and insulating it from the hot surrounding corona. Normal and inverse polarity models of the prominences have been proposed by Kippenhahn & Schlüter (1957) and Kuperus & Raadu (1974) respectively. The relationship between prominence eruption and CMEs has been investigated by many authors (Schmieder *et al.* 2002; Gopalswamy *et al.* 2003; Lin 2004, and the references therein). But the trigger mechanism as well as the overall association between prominence eruption and CME is not well understood. In the present study, we have tried to explore the association between prominence eruption and two CMEs of 21 April 2001.

2. Observations

To carry out the present study, we have used H α images taken from 15 cm coude refractor equipped with CCD photometric camera and Bernhard Halle H α filter from ARIES Nainital, SOHO/EIT (195 Å) and LASCO C2 and C3 white light coronagraph data. The spatial resolution of H α images is 1.3".

3. Analysis

3.1 Morphology and height-time analysis

The prominence eruption as well as the eruption of two associated CMEs occurred during the decay phase of NOAA AR 9433 on 21 April 2001. Since the prominence was not located at the active region no flare or radio burst was found to be associated with this eruption. The spatial correlation between prominence and CMEs can be understood with the help of $H\alpha$ filtergrams, EIT 195 Å difference image and LASCO observations of CMEs (Fig. 1). Comparing the position of prominence eruption with the position angle of CMEs suggests that these events are spatially correlated. To study the temporal relationship between these events the height-time profiles of these events are plotted and compared (Fig. 2).

In Fig. 2 (left) the height-time profiles of the prominence and the leading edge of the first CME are plotted. At 07:50:47 UT the leading edge of the prominence was at $1.02 R_{\text{sun}}$ whereas the extrapolated height of the first CME at 07:50 UT was at $1.0 R_{\text{sun}}$. The plot reveals strong temporal association between these two phenomena. The height-time profile of the second CME is plotted in Fig. 2 (right). The position

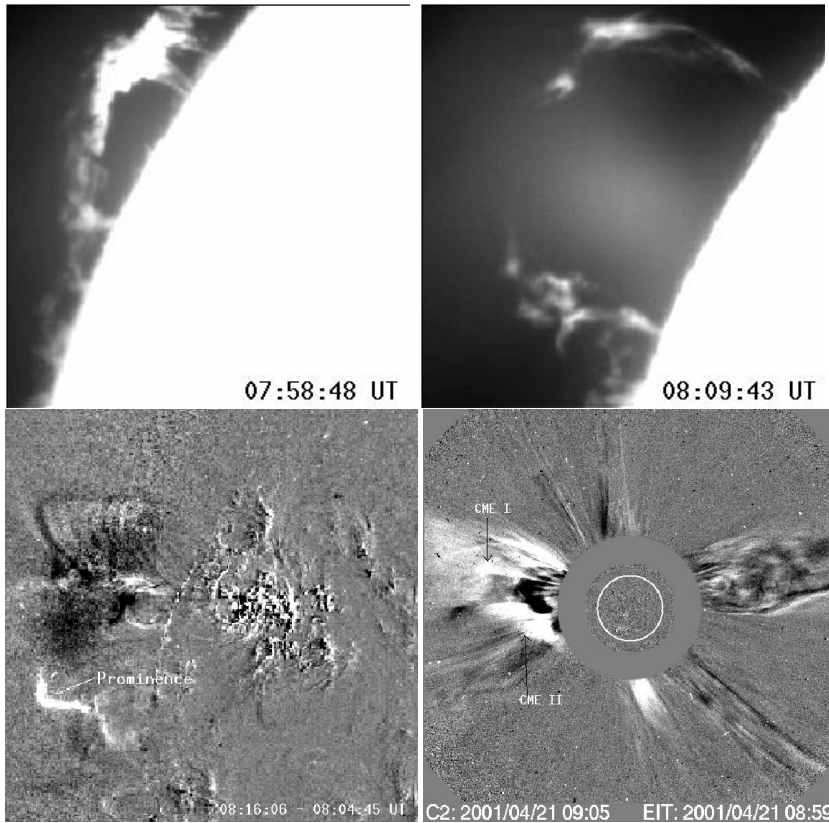


Figure 1. Evolution of $H\alpha$ ($333'' \times 333''$) prominence (**top panel**), prominence eruption in EIT (195 Å) difference image ($910'' \times 910''$) (**bottom left panel**) and associated CMEs (**bottom right panel**). North is up and west is towards right.

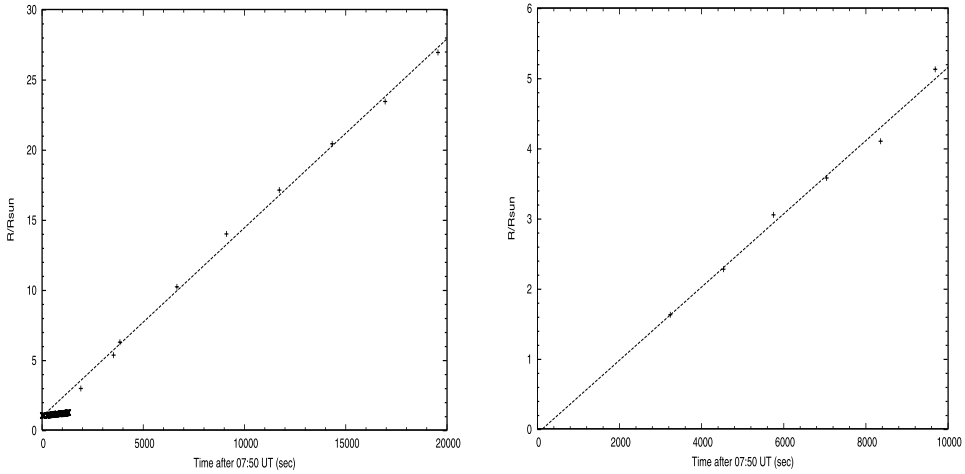


Figure 2. Height-time profiles of first CME (+) and prominence (x) (**left**) and height-time profile of second CME (**right**).

angles of the second CME and the location of SE foot-point suggests that they are spatially correlated with each other. Since the eruption was so huge that it was not possible to keep the entire prominence in the field of view. Thus the rise of SE foot-point could not be covered properly. But whatever information we could gather from our data, suggests that the second CME and burst of SE foot-point are temporally associated with each other within the range of ± 10 minutes. Thus the first CME is spatially and temporally associated with the prominence eruption while the second CME shows good spatial and temporal association with the burst of SE foot-point.

3.2 Energetics

3.2.1 Energy of prominence

The calculation of the energy of prominence is carried out according to the model of Martens (1987). The parameters of prominence are directly derived from $H\alpha$ filter-grams. The momentum equation of an active filament/prominence in Gaussian units is given by (Martens & Kuin 1989),

$$m \frac{d^2 h}{dt^2} = \frac{l I^2}{c^2 h} - \frac{l I B}{c} - \frac{m g_s R_s^2}{(R_s + h)^2}, \quad (1)$$

where R_s is the radius of the Sun, g_s is gravitational acceleration at the solar surface, h is the height above the limb of the Sun, and $B(= B_0 e^{-\frac{\pi h}{D}})$ is the background magnetic field, which is potential in nature. B_0 is the magnetic field in the photosphere and is taken as 50 gauss (Khan *et al.* 1998). D is the half of foot-point separation as measured to be 8.5×10^9 cm. When $(d^2 h)/(dt^2) = 0$, the balance current is derived (Van Tend & Kuperus 1978) for equilibrium height, $h = 1.02 \times 10^{10}$ cm. The loop length of the prominence at this height is measured to be, $l = 2.9 \times 10^{10}$ cm and mass, $m = 0.5 \times 10^{15}$ g. Using these values the strength of equilibrium current is about 0.4×10^{21} stat ampere (s.a.).

The energy stored in the prominence/filament is given as (Martens 1987)

$$E = \frac{1}{2}LI^2 + \frac{1}{c}\phi I, \quad (2)$$

where L is the self inductance of the circuit and ϕ is the magnetic flux of the background field through the circuit. Using the techniques of Wu *et al.* (2002) for the calculation of L and ϕ , the pre-eruptive available energy of the prominence comes out to be 1.2×10^{32} erg.

3.2.2 Energy of CMEs

We examine the energetics of coronal mass ejections (CMEs) with data from LASCO C2 and C3 coronagraph. The enthalpy and thermal energy are so small that they can be

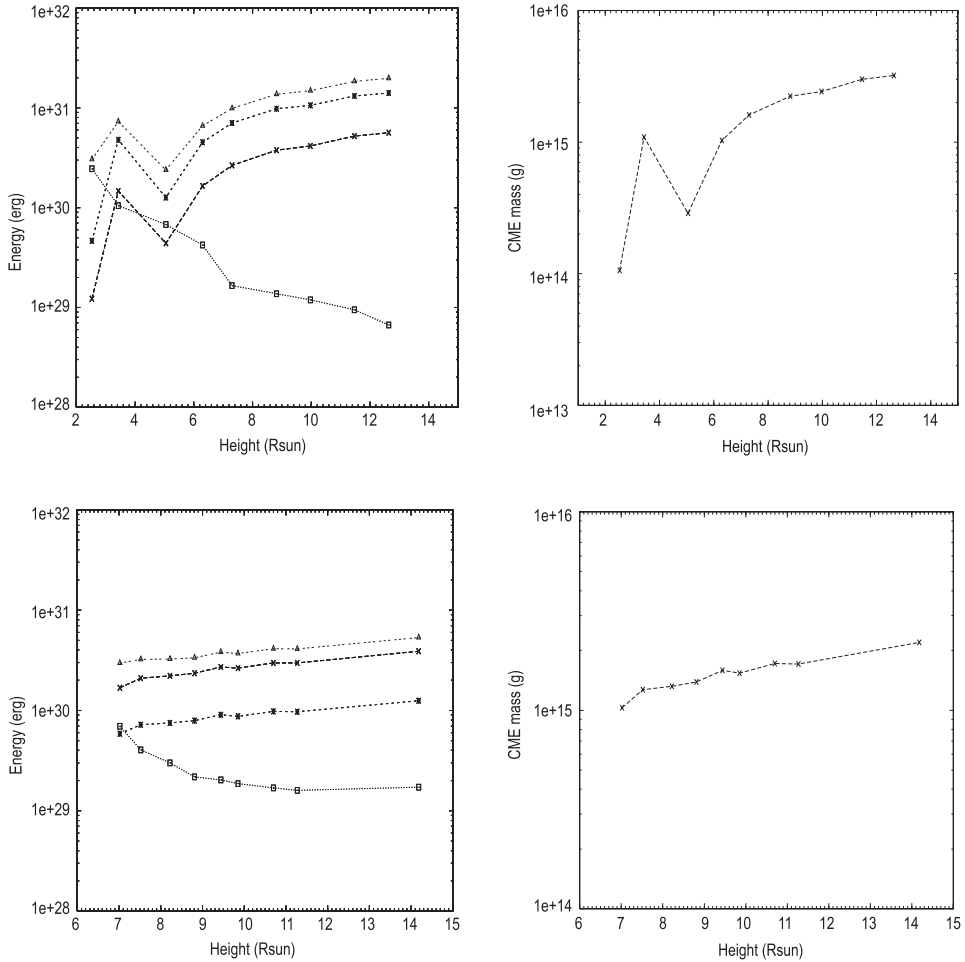


Figure 3. Left panels: Variation of potential (x), kinetic (★), magnetic (□) and total (△) energy of first (top) and second (bottom) CMEs. Right panels: Evolution of mass of first (top) and second (bottom) CMEs.

neglected without effecting the overall conclusions. The potential (E_p), kinetic (E_k) and magnetic (E_m) energies are given by (Vourlidas *et al.* 2000)

$$E_p = GM_s m \left(\frac{1}{R_s} - \frac{1}{R_s + h} \right), \quad (3)$$

$$E_k = \frac{1}{2} m v^2, \quad (4)$$

$$E_m = \frac{l}{8\pi} \times \frac{(B \times A)^2}{A}, \quad (5)$$

where A and l are the area and length of flux rope respectively. l is assumed as $l = r_{\text{cme}}$ (heliocentric height of center of mass). The speeds of the first and second CMEs are 937 km/s and 373 km/s respectively. We have used average magnetic flux $(B \times A) = 1.2 \times 10^{21}$ Max (Wu *et al.* 2002). The variation of potential energy (E_p), kinetic energy (E_k), magnetic energy (E_m) and total energy of both the CMEs with their heliocentric height as well as the evolution of mass of the respective CMEs are shown in Fig. 3. The total energy of the first and second CMEs are 2×10^{31} erg and 5×10^{30} erg respectively.

4. Conclusions

The following conclusions are drawn from the study of prominence eruption and associated CMEs on 21 April 2001.

The spatial and temporal correlation between prominence eruption and initiation of first CME is found to be good. This analysis suggests that both the events are caused by a common disturbance.

The combined energy of both the CMEs is found to be 2.5×10^{31} erg, whereas the energy stored in the prominence is about 10^{32} erg. These values suggest that there was enough energy stored in the prominence to provide both the CMEs.

The total energy of both the CMEs remain constant within a factor of 2, which supports the view that CMEs are magnetically driven and internally powered (Vourlidas *et al.* 2000; Manoharan *et al.* 2001).

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Magnetosonic Waveguide Model of Solar Wind Flow Tubes

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Abstract. We consider solar wind flow tubes as a magnetosonic waveguide. Assuming a symmetric expansion in edges of slab-modelled waveguide, we study the propagation characteristics of magnetosonic wave in the solar wind flow tubes. We present the preliminary results and discuss their implications.

Key words. Solar wind—MHD waves—magnetic flux tubes.

1. Introduction

Parker's assumption (1963) that the solar wind may be fine-structured in the form of flow tubes has also been supported by the HELIOS spacecraft observations (Thieme *et al.* 1990). The wave propagation characteristics in such a magnetically-structured and inhomogeneous medium have been investigated by Roberts (1981a, b). Magnetosonic waves may be trapped in solar wind flow tubes assuming Alfvén speed to be less inside the tube than outside. The presence of the minimum Alfvén speed is required for trapping the waves. This is also correct in a low- β plasma. The trapping of waves can also be caused by the discontinuity of the flow. So, the structures may support the propagation of trapped magnetosonic waves with stable transversal structures determined by the boundary conditions (Nakariakov *et al.* 1996). The trapping of magnetosonic waves are due to reflection of waves from the tangential discontinuity of plasma velocity or due to sudden variation of Alfvén or sound speed. Surface magnetosonic wave is evanescent both inside and outside of waveguide, while the body magnetosonic wave is oscillatory inside the waveguide and evanescent outside. Both the wave modes are localized and non-leaky.

Nakariakov *et al.* (1996) corrected the boundary conditions reported by Mann *et al.* (1992) and have considered the steady flow inside and outside waveguide to study the properties of trapped linear magnetosonic waves and also described the excitations of different modes of magnetosonic waves within the solar wind flow tubes. Joarder *et al.* (1996) have also studied the magnetosonic modes of magnetic structures in the presence of inhomogeneous steady flows. Joarder and Narayanan (2000) have examined the combined effect of non-parallel propagation and steady shear flows on the properties of hydromagnetic surface waves. Considering the non-parallel propagation of magnetosonic waves, Joarder (2002) have investigated these waves in the observed fine structures of high-speed solar wind streams. In this paper, we study the propagation

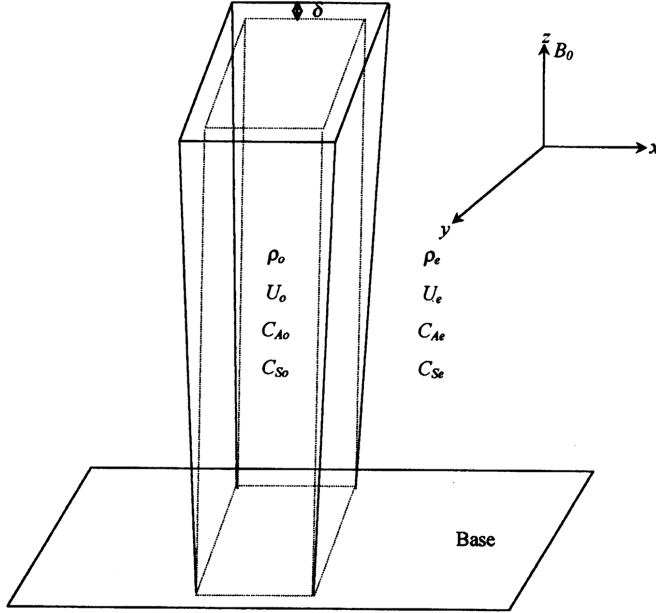


Figure 1. The solar wind flow tube model.

properties of magnetosonic waves. We also consider a symmetric expansion (δ) in the edges of the slab and its effect on propagation properties. We have studied the effect of magnetic field geometry of solar wind flow tubes on the propagation properties of different trapped modes of magnetosonic waves. Our solar wind flow tube model is more general in the sense that it supports both the body and surface magnetosonic waves. Each of the type of waves is also classified as kink and sausage modes.

2. Waveguide model and dispersion relation

We consider a magnetic slab with finite dimension $x = \pm d$ and $y = \pm d$, having a background magnetic field in the z -direction, assuming the slab to be infinitely long. The steady shear flow along the z -axis are U_0 and U_e for inside and outside the medium respectively. The sound speeds are C_{S0} and C_{Se} for inside and outside the slab respectively and similarly Alfvén speeds are C_{A0} and C_{Ae} . Plasma densities are ρ_0 and ρ_e inside and outside the slab respectively.

We consider a symmetric expansion in the edges of the slab (δ) (Fig. 1). This follows the flux conservation law

$$B_{0z} A_{0z} = B_z A_z, \quad (1)$$

where B_{0z} and B_z are magnetic fields at the base and at any arbitrary height of the slab respectively. The respective cross-sectional areas of the slab are A_{0z} and A_z .

Using equation (1) and the slab geometry shown in Fig. 1, we get

$$\delta = \pm \frac{B_{0z} d^2}{B_z} - \frac{d^2}{4d}. \quad (2)$$

Using the dimensionless parameter $d \approx 1$, we have

$$\delta = \pm \frac{B_{0z} - B_z}{4B_z}. \quad (3)$$

δ represents an areal expansion of the slab at any height which may affect the transversal velocity structure both inside and outside the slab and can be incorporated in the coordinates. The positive sign in equations (2, 3) corresponds to divergent magnetic field lines geometry while the negative sign corresponds to convergent magnetic field lines geometry in our solar wind flow tubes model. We used only $+\delta$ in our calculations, i.e., divergent magnetic field lines geometry of solar wind flow tubes. We have studied the effect of this magnetic field geometry on the propagation properties of magnetosonic wave modes.

The linearized MHD equations for ideal MHD give the following differential equation for inside and outside the slab (cf., Mann *et al.* 1992; Nakariakov *et al.* 1996).

$$\frac{d^2 V_{xi}}{dx^2} - m_i^2 V_{xi} = 0, \quad (4)$$

where i is either for inside or for outside the slab. The transversal plasma velocity is $V_{xi}(x) \exp i(\omega t - kz)$, and

$$m_i^2 = \frac{[a_{Ai}^2 - (a - M_i)^2][a_{Si}^2 - (a - M_i)^2]}{a_{fi}^2[a_{Ti}^2 - (a - M_i)^2]} k^2, \quad (5)$$

where,

$$\begin{aligned} a &= \frac{\omega}{kC_{A0}}; & a_{Ai}^2 &= \frac{C_{Ai}^2}{C_{A0}^2}; & a_{Si}^2 &= \frac{C_{Si}^2}{C_{A0}^2}; \\ M_i &= \frac{U_i}{C_{A0}}; & a_{Ti}^2 &= \frac{C_{Ai}^2 C_{Si}^2}{C_{A0}^2 (C_{Ai}^2 + C_{Si}^2)}; & a_{fi}^2 &= \frac{C_{Ai}^2 + C_{Si}^2}{C_{A0}^2}. \end{aligned}$$

All variables are normalized with C_{A0} , M_i is Alfvén Mach number and a is phase speed in units of Alfvén speed C_{A0} . a is the phase speed of different magnetosonic modes. Boundary conditions are as follows (cf., Nakariakov *et al.* 1996):

$$\frac{V_{x0}(x = \pm d)}{\omega - kU_0} = \frac{V_{xe}(x = \pm d)}{\omega - kU_e}, \quad (6a)$$

$$p_{To}(x = \pm d) = p_{Te}(x = \pm d) \quad (6b)$$

and

$$p_{Ti}(x) = \frac{iC_{A0}\rho_i a_{fi}^2 [a_{Ti}^2 - (a - M_i)^2]}{k(a - M_i)[a_{Si}^2 - (a - M_i)^2]} \frac{dV_{xi}}{dx}. \quad (6c)$$

The solutions for the transversal structures of modes outside the slab in our model are given by,

$$V_{xe}(x) = \left\{ \begin{array}{ll} A_1 \exp[-m_e(x \pm \delta - d)], & x > d \\ A_2 \exp[+m_e(x \pm \delta + d)], & x < d \end{array} \right\}, \quad (7)$$

where A_1 and A_2 are constants. This equation is valid only in the case when the wavelength of the perturbation is much shorter than the characteristic spatial scale of the tube expansion according to the WKB approximation.

Solutions inside the slab for surface and body waves are:

$$V_{x0}(x) = \left\{ \begin{array}{ll} A \sinh[m_0(x \pm \delta)], & \text{for sausage surface modes} \\ A \cosh[m_0(x \pm \delta)], & \text{for kink surface modes} \\ A \sin[n_0(x \pm \delta)], & \text{for sausage body modes} \\ A \cos[n_0(x \pm \delta)], & \text{for kink body modes} \end{array} \right\}, \quad (8)$$

where A is constant and $n_0^2 = -m_0^2$.

We have taken $M_e = 0$ and $M_0 \rightarrow M$ because we have considered in our calculations the frame of reference related to the external flow. The M value is given by $M = (U_0 - U_e)/C_{Ao}$ which represents the difference between Alfvén Mach number inside and outside of solar wind flow tubes. Some unspecified acceleration mechanism is likely to be present inside flow tubes, so we have taken the constant value of steady flow. Nakariakov *et al.* (1996) also used the same approach in his calculations related to excitations of body magnetosonic waves in solar wind flow tubes. Applying the boundary conditions (6) on solutions given by equations (7–8), we also consider the incompressible plasma ($\gamma \rightarrow \infty$), for which m_0 (or n_0) and m_e tends to k . We get the following dispersion relations for surface magnetosonic waves:

$$\left[\tanh(kd \pm k\delta) + \frac{\rho_e}{\rho_0} \right] a^2 - 2Ma \tanh(kd \pm k\delta) + (M^2 - 1) \tanh(kd \pm k\delta) - \frac{\rho_e}{\rho_0} a_{Ae}^2 = 0, \text{ for kink surface wave} \quad (9)$$

and

$$\left[\coth(kd \pm k\delta) + \frac{\rho_e}{\rho_0} \right] a^2 - 2Ma \coth(kd \pm k\delta) + (M^2 - 1) \coth(kd \pm k\delta) - \frac{\rho_e}{\rho_0} a_{Ae}^2 = 0, \text{ for sausage surface wave.} \quad (10)$$

When plasma β is small, then the Alfvén wave exceeds the sound speed. Surface mode may or may not exist for low- β plasma. There cannot be two-sided surface waves in a low- β plasma and it will not hold if only one side of the magnetic interface has a low- β (Roberts 1981a; Wentzel 1979). That means the existence of one-sided normal modes of surface waves. Under this limit following two relations given for single normal mode of body waves, which may excite due to tangential discontinuity.

$$\left[\tan(kd \pm k\delta) \pm \frac{\rho_e}{\rho_0} \right] a^2 - 2Ma \tan(kd \pm k\delta) + (M^2 - 1) \tan(kd \pm k\delta) - \frac{\rho_e}{\rho_0} a_{Ae}^2 = 0, \text{ for kink body waves} \quad (11)$$

and

$$\left[\frac{\rho_e}{\rho_0} - \cot(kd \pm k\delta) \right] a^2 + 2Ma \cot(kd \pm k\delta) + (1 - M^2) \cot(kd \pm k\delta) - \frac{\rho_e}{\rho_0} a_{Ae}^2 = 0, \quad \text{for sausage body waves.} \quad (12)$$

Dimensionless parameter $(kd \pm k\delta)$ in degree ($^\circ$), is employed in equations (9–12).

3. Magnetic slab with flow

We have considered a flow ($M = 1.46$) and actual solar-wind data of Thieme *et al.* (1990). We have used the parameters of solar wind flow tubes, Alfvén speed inside solar wind flow tubes $C_{Ao} = 65 \text{ km s}^{-1}$, sound speed inside solar wind flow tubes $C_{So} = 65 \text{ km s}^{-1}$, outside of flow tubes Alfvén speed $C_{Ae} = 100 \text{ km s}^{-1}$ and sound speed $C_{Se} = 70 \text{ km s}^{-1}$. Steady flow inside of flow tubes $U_0 = 750 \text{ km s}^{-1}$ and outside it $U_e = 655 \text{ km s}^{-1}$. $\rho_e/\rho_0 = 0.583$, $a_{Ae}^2 = 2.37$, $d = 1.6 \times 10^6 \text{ km}$ are also used in our calculations. Figure 2 shows increasing phase velocity pattern for kink-surface modes. The phase velocities are greater for expanded tubes compared to the uniform magnetic flux-tubes. However, the flux-tube geometry does not have any effect at lower wavelengths.

Figure 3 shows two sausage surface modes. This could possibly be categorized as fast sausage surface wave and slow sausage surface wave. The lower limit of the fast sausage surface wave is $(a_{Ao} + M)$. The upper limit of slow sausage surface wave is found at $(M - a_{Se})$. The effect of boundary perturbations is decreased for the phase velocities of two kinds of surface-sausage waves, but not affected at lower wavelengths.

Figures 4 and 5 show the phase velocity pattern of kink and sausage body waves respectively. Both the modes exhibit the increasing phase velocity pattern with velocities of two types of modes. We have plotted the phase speeds (a) of kink and sausage

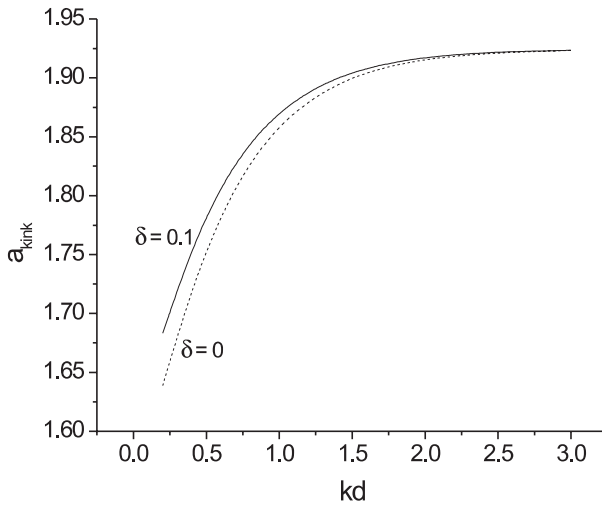


Figure 2. Dispersion curves for kink mode of surface magnetosonic wave in magnetic slab with flow ($M = 1.46$).

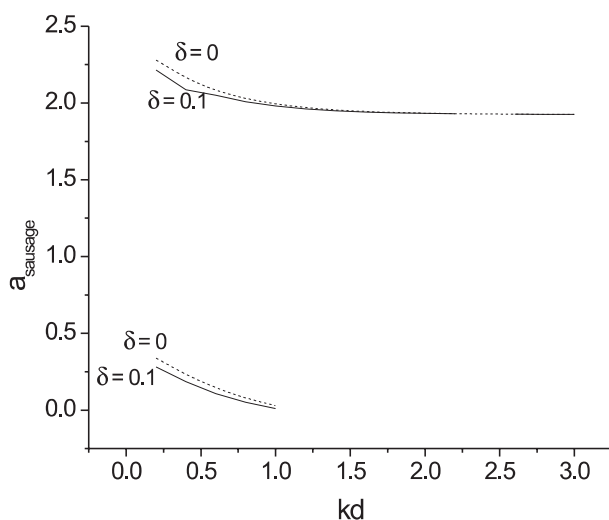


Figure 3. Dispersion curves for sausage mode of surface magnetosonic wave in magnetic slab with flow ($M = 1.46$).

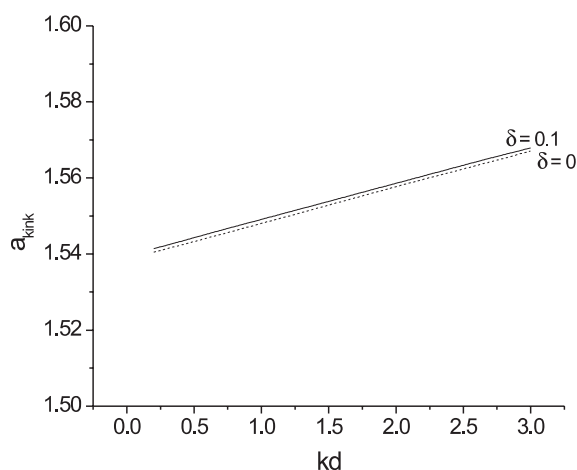


Figure 4. Dispersion curves for kink mode of body magnetosonic wave in magnetic slab with flow ($M = 1.46$).

modes of surface and body magnetosonic waves with respect to kd in Figs. 2–5. kd is a dimensionless number. Since waveguide approach supports waves having wavelengths of the order of d or greater than d (i.e., $\lambda \geq d$), so the wave number k multiplied by d yields a dimensionless number.

4. Concluding remarks

We have studied the propagation characteristics of different modes of magnetosonic waves, considering a magnetic slab with different outside and inside steady flows. We

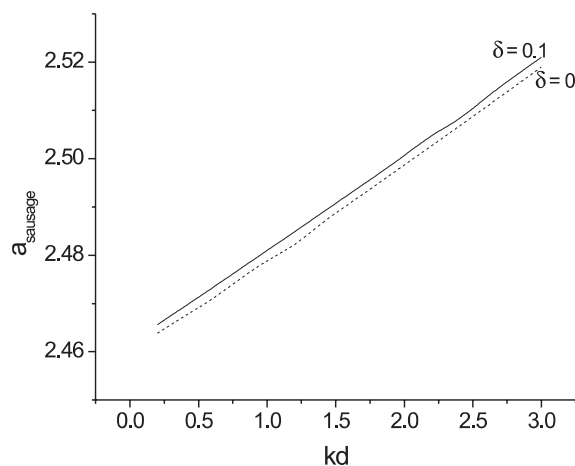


Figure 5. Dispersion curves for sausage mode of body magnetosonic wave in magnetic slab with flow ($M = 1.46$).

also considered a symmetric expansion in edges of the slab at any arbitrary height. We find that surface waves may exist due to magnetic field discontinuity. When we approximated the tube embedded in plasma with different plasma- β inside the tube, we see that it supports the existence of surface waves. We also find one normal and single mode of body wave due to tangential discontinuity and we have not considered higher modes of compressible plasma. With steady flow ($M = 1.46$), kink mode velocity increases as kd increases while sausage mode of surface wave exhibits decreasing velocity pattern with kd . In case of sausage modes with flow, two modes exist. The first has a lower limit of phase speed ($a_{A0} + M$) while the second has an upper limit of phase speed ($M - a_{se}$). These may be categorized as fast and slow modes. Kink and sausage body waves show increasing velocity pattern. The ' δ ' factor does not affect the phase velocities, particularly at higher wave number.

Acknowledgements

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Analysis of Plasma and Field Conditions during Some Intensely Geo-effective Transient Solar/Interplanetary Disturbances of Solar Cycle 23

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Abstract. The problem of solar wind–magnetosphere coupling is investigated for intense geomagnetic storms ($Dst \leq -100$ nT) that occurred during solar cycle 23. For this purpose interplanetary plasma and field data during some intensely geo-effective transient solar/interplanetary disturbances have been analysed. A geomagnetic index that represents the intensity of planetary magnetic activity at subauroral latitude and the other that measures the ring current magnetic field, together with solar plasma and field parameters (V , B , B_z , σ_B , N , and T) and their various derivatives (BV , $-BV_z$, BV^2 , $-BzV^2$, B^2V , Bz^2V , NV^2) have been analysed in an attempt to study mechanism and the cause of geo-effectiveness of interplanetary manifestations of transient solar events. Several functions of solar wind plasma and field parameters are tested for their ability to predict the magnitude of geomagnetic storm.

Key words. Earth: geomagnetic storm—Sun: solar and wind—magnetosphere coupling.

1. Introduction

Magnetic storms of various intensities and their causes were studied by various research groups (e.g., Burlaga *et al.* 1987; Tsurutani *et al.* 1992; Badruddin 1998; Dubey and Mishra 2001; Echer *et al.* 2004; Zhang *et al.* 2004), and there has been substantial growth in our knowledge of solar and interplanetary causes of geomagnetic storms (Crooker 2000; Gopalaswamy 2004). However, there are still unanswered questions that must be addressed and solved to predict the occurrence of magnetic storms (e.g., see Tsurutani and Gonzalez 1997). In order to understand the response of the magnetosphere to interplanetary conditions during magnetic storms, several studies have derived relations between interplanetary values of solar wind velocity, magnetic field, its north–south component and various combinations of these parameters but a unique relationship which may ultimately lead to an unambiguous understanding of the phenomena and predict their occurrence, is still unknown.

The purpose of this paper is to study the solar wind–magnetosphere coupling problem during some intensely geo-effective transient solar/interplanetary disturbances of solar cycle 23.

2. Results and discussion

We have selected nine intensely geo-effective storms ($Dst \leq -100$ nT) of solar cycle 23 during which hourly solar plasma and field data were available. One of these storms, plotted in Fig. 1, shows features such as a storm in sudden commencement (SSC), two-step increase in geomagnetic activity, extended period of Dst minimum (~ 20 h) and two-step (a fast and then a slow) recovery. To find the solar source of the storm, we have looked at the table of earth-directed coronal mass ejections (CMEs) (Gopalaswamy *et al.* 2001; Cane and Richardson 2003; Manoharan *et al.* 2004). This storm is due to an interplanetary CME (ICME) of August 25, 1998 (Cane and Richardson 2003). In Fig. 1, we have shown the time profile of the storm (shown by three geomagnetic indices ap , Kp and Dst) along with the simultaneous variations in solar wind velocity (V), IMF strength (B), its variance (σ_B) and north-south component (B_z), plasma density (N) and its temperature (T). Various functions of these parameters, dynamic pressure (NV^2) and some other parameters BV , BzV , BV^2 , BzV^2 , B^2V and Bz^2V have also been plotted in this figure.

At the time of a sudden positive jump in Dst (an SSC), there is a sharp increase in velocity (V) and density (N), hence dynamic pressure NV^2 coincident with arrival of shock at 1800 hours on August 25, 1998. This SSC is a result of sudden compression of the earth's magnetosphere. After SSC, within four hours, B_z becomes negative and geomagnetic activity starts increasing, reaching to a maximum level in two steps, remains high (~ -150 nT) for an extended period (~ 20 hours) ($B_z \leq -10$ nT for ~ 20 hours) and the electric field is > 5 mV/m for about 24 hours during this period and

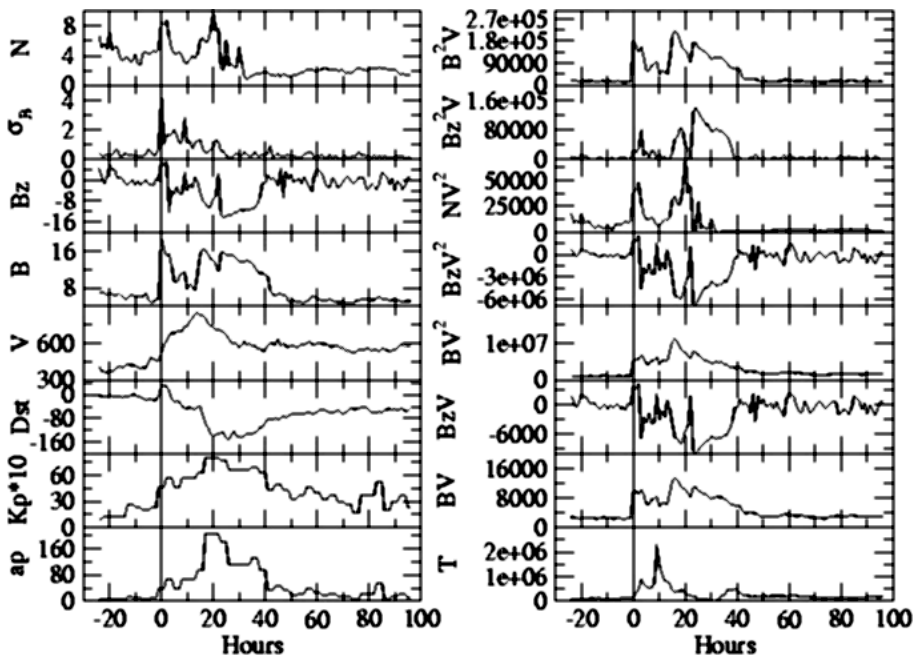


Figure 1. Variation of geomagnetic parameters (ap , Kp , Dst), solar wind plasma and field parameters (V , B , B_z , σ_B , N , T) and their various functions due to a shock associated CME event of August 25, 1998 (zero hour corresponds to the arrival time of the disturbance).

then recovery takes place in two steps, a sharp recovery from -111 nT to -72 nT in just 9 hours and then a much slower recovery towards pre-decrease level. A look at the time profile of the geomagnetic indices, interplanetary plasma and field parameters, and their various functions, plotted in Fig. 1, shows simultaneous variations in a number of panels, however, negative excursion in B_z values appears to be better related to the initial phase of the storm. For about 15 hours (typical time for sheath passage), after a few (five) hours of the arrival of shock, the B_z is largely negative but fluctuating, indicating that this field is caused by the amplification of the pre-existing southward/northward field in the sheath region by shock compression. This region is followed by a magnetically quiet region (low σ_B) of large southward field of a long duration.

Another event (Fig. 2) shows a one-step geomagnetic storm that occurred on April 06, 2000 at 1600 hours. This storm was not preceded by an SSC probably due to small change in NV^2 at the shock arrival. However, a large geomagnetic storm

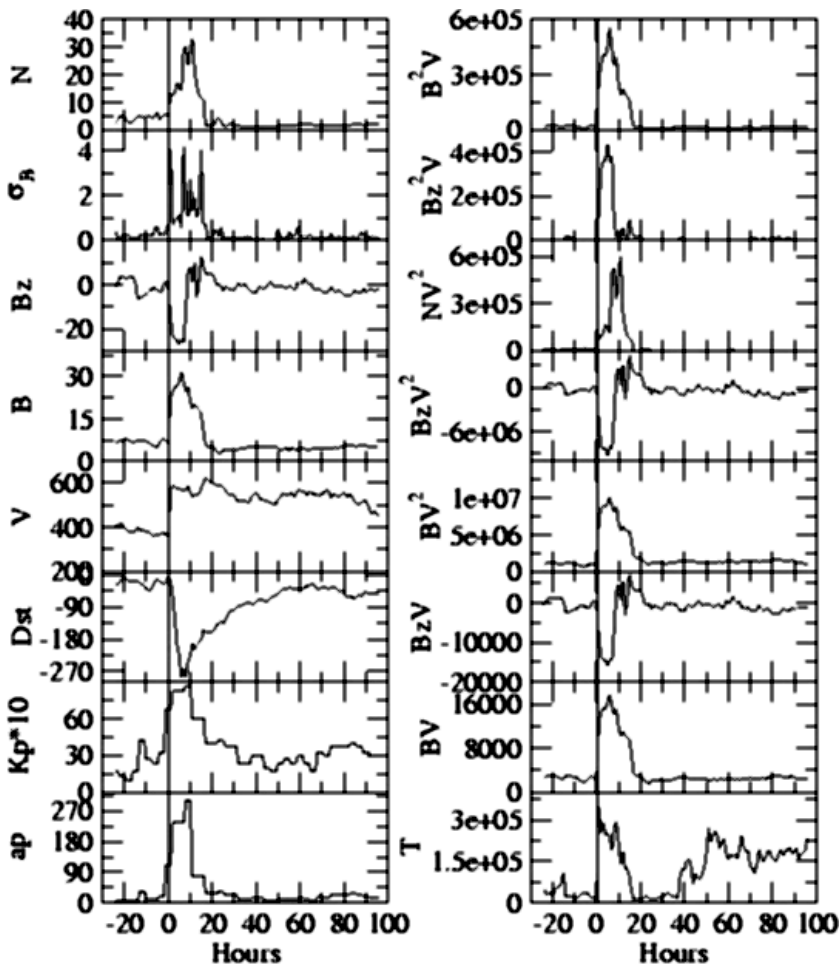


Figure 2. Variation of geomagnetic parameters (ap , Kp , Dst), solar wind plasma and field parameters (V , B , B_z , σ_B , N , T) and their various functions due to a shock associated CME event of April 06, 2000 (zero hour denotes the arrival time of the disturbance).

Table 1. Start time and peak values of geomagnetic and interplanetary parameters during nine selected geomagnetic storm of solar cycle 23.

Event	Y/M/D	Time	Dst	ap	Kp	V	B	σ_B	Bz
1	98/08/25	1800	-154	200	8.0	838	19	4.1	-24
2	00/04/06	1600	-284	292	8.7	621	31	4.2	-26
3	01/03/30	2300	-384	299	8.7	827	50	23.3	-44
4	01/04/11	1200	-258	234	8.4	827	34	5.2	-20
5	02/04/17	1300	-128	132	7.0	608	31	7.7	-134
6	02/04/19	0700	-143	154	7.3	664	22	5.2	-15
7	03/11/20	0800	-472	300	8.8	698	55	5.0	-50
8	04/07/24	0400	-150	208	8.0	596	22	4.9	-18
9	04/07/26	0200	-174	272	8.7	975	26	4.5	-19

Table 1. (Continued.)

Event	$BV \times 10^4$	$BzV \times 10^3$	$B^2V \times 10^5$	$Bz^2V \times 10^6$	$BV^2 \times 10^7$	$BzV^2 \times 10^5$	$NV^2 \times 10^4$
1	1.34	-9.05	2.20	0.14	1.13	-1.4	6.88
2	1.75	-14.80	5.59	0.38	1.00	-83.0	52.42
3	3.24	-30.26	15.00	1.37	2.10	-209.7	102.30
4	2.54	-13.79	8.74	2.92	1.82	-105.3	42.55
5	1.46	-8.69	4.49	0.17	0.73	-33.8	47.55
6	1.47	-7.66	3.21	0.12	0.97	-36.3	8.74
7	33.75	-29.69	19.00	1.59	2.11	-192.5	23.57
8	1.31	-10.38	2.83	0.19	0.75	-57.0	1.46
9	2.55	-16.99	6.70	0.36	2.52	-150.0	0.85

(Dst ~ -284 nT) results with its peak during a period when Bz is large (and negative) together with enhanced V and B. Further, peak Dst, Kp and ap occur during the passage of sheath region as evident from the enhanced values of V, B, σ_B , N and T. The storm shown in Fig. 2 is a one-step storm while that shown in Fig. 1 is a two-step geomagnetic storm. A table of selected events observed during 1998–2004, with their associated interplanetary parameters is given in Table 1. The results shown in Figs. 1 and 2 (see also Table 1) concur with the suggestion that the primary cause of intense geomagnetic storm is the occurrence of intense ($Bz \leq -10$ nT) and long duration southward IMF, and in such conditions, the energy transforms from solar wind to the magnetosphere more efficiently due to merging/reconnection mechanism.

To be able to forecast the magnitude of an impending geomagnetic storm is a major goal of space weather research. We have performed a regression analysis (Fig. 3) of peak values ($-Dst/ap$) of nine intense geomagnetic storms versus the peak values of solar wind plasma/field parameters (V, B, Bz) and their various functions (BV, $-BzV$, B^2V , $-BzV^2$, BV^2 and Bz^2V). The linear correlation coefficients so obtained are given in Table 2. Although, the correlation coefficient is maximum for duskward electric field ($-BzV$); however, given the uncertainties in their values shown in Table 2, the significance of correlation between Dst and parameters such as B, Bz, VB^2 cannot be treated as less significant. Similarly, although the correlation between ap and $-BzV$

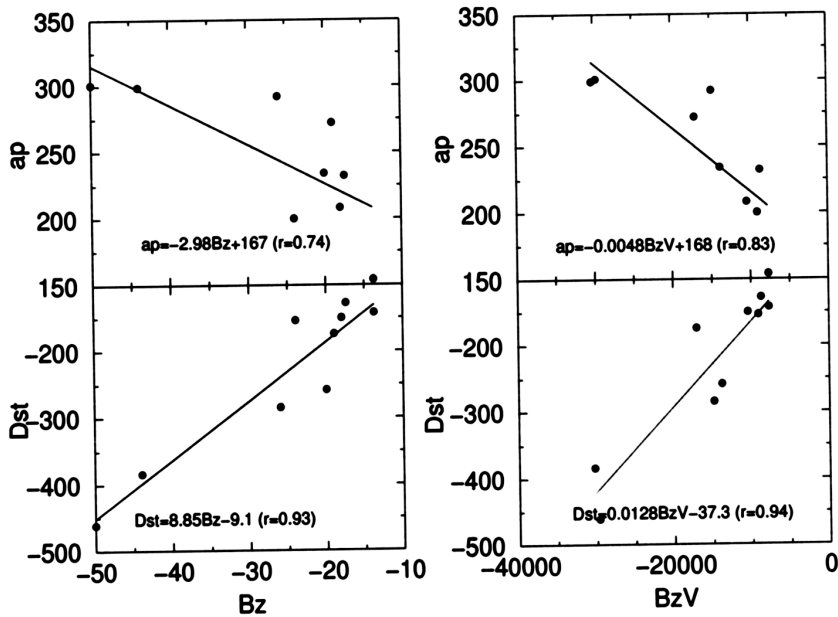


Figure 3. Scatter plot between peak values of geomagnetic parameters (Dst and ap) and interplanetary parameters (southward component of the field ($-B_z$) and dawn-dusk electric field ($-B_zV$)) for nine intense geomagnetic storms.

Table 2. Correlation coefficient with probable error between various interplanetary (IP) and geomagnetic parameters (GP).

IP/GP	Dst	ap
V	-0.08 ± 0.002	0.34 ± 0.108
B	-0.92 ± 0.035	0.67 ± 0.123
B_z	0.93 ± 0.030	-0.74 ± 0.102
BV	-0.74 ± 0.101	-0.45 ± 0.179
B_zV	0.94 ± 0.026	-0.83 ± 0.070
B^2V	-0.91 ± 0.039	0.72 ± 0.108
B_z^2V	-0.01 ± 0.224	0.03 ± 0.224
BV^2	-0.56 ± 0.154	0.73 ± 0.105
B_zV^2	0.26 ± 0.209	0.33 ± 0.200

is highest, the correlation with other parameters such as B_z , VB^2 and BV^2 is sufficiently large. It is also clear from the table that the correlation coefficient of solar wind parameters is, in general, smaller with ap than Dst index. One possible reason for this difference may be that the ap values are 3-hourly values while Dst and solar parameters are hourly data. The results also suggest the need to perform analysis with a larger number of such events. We hope to use a larger database in future analyses.

The best-fit linear equation between the magnitude of geomagnetic storms and peak values of VBz is obtained:

$$ap_{\text{peak}} = -0.00483 \times (VBz)_{\text{peak}} + 167.63,$$

$$Dst_{\text{peak}} = 0.012739 \times (VBz)_{\text{peak}} - 37.34,$$

where Bz is in nT and V in km/sec.

These equations may be used to obtain the magnitude of storm if the maximum velocity of solar/interplanetary disturbance (see also, Srivastava and Venkatakrishnan 2002) along with the peak value of the north–south component of magnetic field during a disturbance is known.

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The Influence of Solar Activity on the Rainfall over India: Cycle-to-Cycle Variations

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Abstract. We use 130 years data for studying correlative effects due to solar cycle and activity phenomena on the occurrence of rainfall over India. For the period of different solar cycles, we compute the correlation coefficients and significance of correlation coefficients for the seasonal months of Jan–Feb (JF), Mar–May (MAM), June–Sept (JJAS) and Oct–Dec (OND) and, *annual mean* data. We find that: (i) with a moderate-to-high significance, Indian rainfall is correlated with the sunspot activity and, (ii) there is an overall trend that during the period of low sunspot activity, occurrence of rainfall is high compared to the period of high sunspot activity.

We speculate in this study a possible physical connection between the occurrence of the rainfall and the sunspot activities and, the flux of galactic cosmic rays. Some of the negative correlations between the occurrences of the sunspot and rainfall activities obtained for different solar cycle periods are interpreted as effects of aerosols on the rain forming clouds due to either intermittent volcanic eruptions or due to intrusion of interstellar dust particles in the Earth's atmosphere.

Key words. Solar activity—Indian rainfall—galactic cosmic rays—volcanic or interstellar dust.

1. Introduction

Being predominantly an agricultural country, the Indian economy depends mainly on rainfall. For the last few years, drought has hit the world's 12th largest economy. The dominant contribution ($\sim 25\%$ of the gross domestic product) towards the country's economy is agriculture which depends mainly upon the monsoon rainfall. For the last successive years, the Indian subcontinent witnessed the miserable failure of southwest monsoon that led to severe droughts. Causes for the abnormal activities of the rainfall over India such as occurrence of floods and droughts is not yet completely understood. It is believed that the main causes are due to the localized anthropogenic influences over the climate and environment, *viz.*, degradation of forest coverage and other unknown causes. Another cause which is clearly discernible in most of the recent findings is due to solar radiative forcing over the global climate and the environment of the Earth. Recent studies (Hiremath and Mandi 2004 and references therein) indicate that the solar cycle and related activity phenomena have a good correlation with the Earth's

global climate and temperature. In the previous study (Hiremath and Mandi 2004), we investigated the long term (> 20 yrs) influence of the solar activity on the Indian monsoon rainfall. In the present study we investigate the effect of the solar activity for a short term (~ 11 yrs) scale and study the cycle-to-cycle variations of the influence of the solar activity on the rainfall over India.

2. Data and analysis

We consider 130 years (1871–2000) data of the sunspot numbers and the rainfall occurrence activity (Parthasarathy *et al.* 1993; <http://www.tropmet.res.in>) over India. The sunspot number data is considered from the National Geophysical Data Center, Boulder, Colorado, USA (<http://www.ngdc.noaa.gov/STP/SOLAR/SSN/ssn.html>). Parthasarathy *et al.* (1993) have compiled a homogeneous set of rainfall data from the 14 meteorological subdivisions covering the northwestern and central parts of India (about 55% of the total area of the country). For the years 1871–2000, the rainfall data (in mm) is available in monthly, seasonal and annual series. In the present analysis, we use the seasonal (including the spring, the southwest, the northeast and the winter monsoon) and the annual (averaged for the period of 12 months) rainfall data.

The different nomenclatures for the seasonal rainfall occurrences are given as follows. We combine the winter rainfall data for the months of January and February and denote as JF. For the combined months of the March, April and May, *spring* rainfall data set, we denote as MAM. In the rest of the months, the rainfall variability is dominated by the southwest (June, July, August and September) and northeast (October, November and December) monsoons which we denote as JJAS and OND respectively. For a particular year, we compute the average of all the 12 months rainfall data and call it as *annual mean* of the rainfall data. For similar combined months (JF, MAM, JJAS and OND), we collect the sunspot data. For each year, by using 12 months data, we compute *annual mean* of the sunspots. In Table 1, we present the the Spearman Rank-Order correlation coefficients (and significances) computed between the occurrences of sunspot and rainfall activities from the raw data. This method (Press *et al.* 1992) of finding the correlation between two variabilities is more robust than the usual method (*i.e.*, by linear correlation). The first column represents the solar cycle period, the second column, the length of the solar cycle, the third column, total area of the rainfall occurrence variability and, the fourth column, the area of the solar cycle. Following Hiremath and Mandi (2004), the length of the solar cycle, area of the solar cycle and area of the rainfall occurrence variability are computed. In the fifth to last columns, the results for the combined seasonal months of JF (winter), MAM (summer), JJAS (southwest), OND (southeast) and the annual mean correlation coefficients have been presented. The significance of correlation coefficients are given in brackets just below the values of correlation coefficients.

The absolute values of the correlation coefficients (seasonal and annual) presented in Table 1 are plotted with respect to the annual area of the solar cycle in the first plot of the Fig. 1. In the rest of the four plots of Fig. 1, we illustrate separately each of the seasonal months' correlation coefficients with the seasonal sunspot cycle area. In the last plot of Fig. 1, correlation coefficients for the annual with respect to the annual solar cycle areas are presented. By taking significance of correlation coefficients as errors in the obtained correlation coefficients, we subject both the seasonal correlation coefficients and solar cycle area data set to the linear least square fit of the form

Table 1. Cycle-to-cycle variation of correlation coefficients for the seasonal months (JF, MAM, JJAS and OND) and annual raw data respectively. The values in brackets are significance of the correlation coefficients. Length of the solar cycle is computed from min-min years.

Year	Length	Rainfall area	Cycle area	JF	MAM	JJAS	OND	Annual
1878–1889	12	11008.5	4633.6	0.09 (0.22)	−0.22 (0.50)	0.13 (0.32)	−0.11 (0.27)	0.08 (0.12)
1890–1901	12	10332.0	5541.0	0.25 (0.56)	−0.10 (0.24)	0.50 (0.91)	0.70 (0.99)	0.55 (0.92)
1902–1913	12	9572.9	4469.8	0.22 (0.51)	0.36 (0.75)	−0.12 (0.29)	−0.41 (0.82)	−0.04 (0.10)
1914–1923	10	7096.8	4459.0	0.16 (0.35)	0.50 (0.86)	0.55 (0.90)	0.55 (0.90)	0.21 (0.44)
1924–1933	10	7355.0	4103.6	0.38 (0.72)	−0.36 (0.69)	−0.20 (0.42)	−0.35 (0.67)	−0.41 (0.76)
1934–1944	11	8941.0	6698.6	−0.16 (0.37)	0.17 (0.39)	0.11 (0.26)	0.38 (0.75)	−0.34 (0.69)
1945–1954	10	7379.4	7514.4	−0.02 (0.04)	0.26 (0.53)	−0.30 (0.60)	0.52 (0.87)	0.22 (0.47)
1955–1964	10	8009.2	9547.0	0.39 (0.78)	0.38 (0.72)	−0.36 (0.69)	0.29 (0.56)	0.07 (0.15)
1965–1976	12	9942.6	8494.7	0.14 (0.34)	0.32 (0.68)	−0.04 (0.10)	−0.27 (0.61)	−0.05 (0.12)
1977–1986	10	7137.7	8306.3	0.03 (0.07)	−0.03 (0.07)	0.16 (0.35)	−0.13 (0.27)	0.30 (0.60)
1987–1996	10	7149.9	7802.0	−0.73 (0.99)	0.25 (0.50)	0.36 (0.69)	−0.46 (0.81)	−0.01 (0.01)

$R = a + bS$ (R is the correlation coefficient, S is the area of the solar activity, a and b are usual constants). The results obtained from such a linear relation with goodness of fit is inserted at the top of each plot. The red continuous line over plotted in each of the plot is computed from the least square fit.

One can argue that the correlations (presented in Table 1) within a cycle are rarely significant to draw any conclusion. The following are the two most likely reasons for the low significance of correlation coefficients (CC). The first problem is estimation of the significance of correlation coefficients (CC). The best and correct way is to have measured uncertainties in the rainfall data and compute the CC and their significances. Unfortunately we don't have any measured uncertainties in the rainfall data considered in this study. Thus we have considered $(N)^{1/2}$ (where N is the number data points in each variable) as measured uncertainty in each rainfall data. The second problem is the low accuracy of the rainfall data that are measured from the rain gauges.

Though the uncertainties in the derived coefficients are large, we can draw the following conclusions: for a particular solar cycle period, there appears to be an

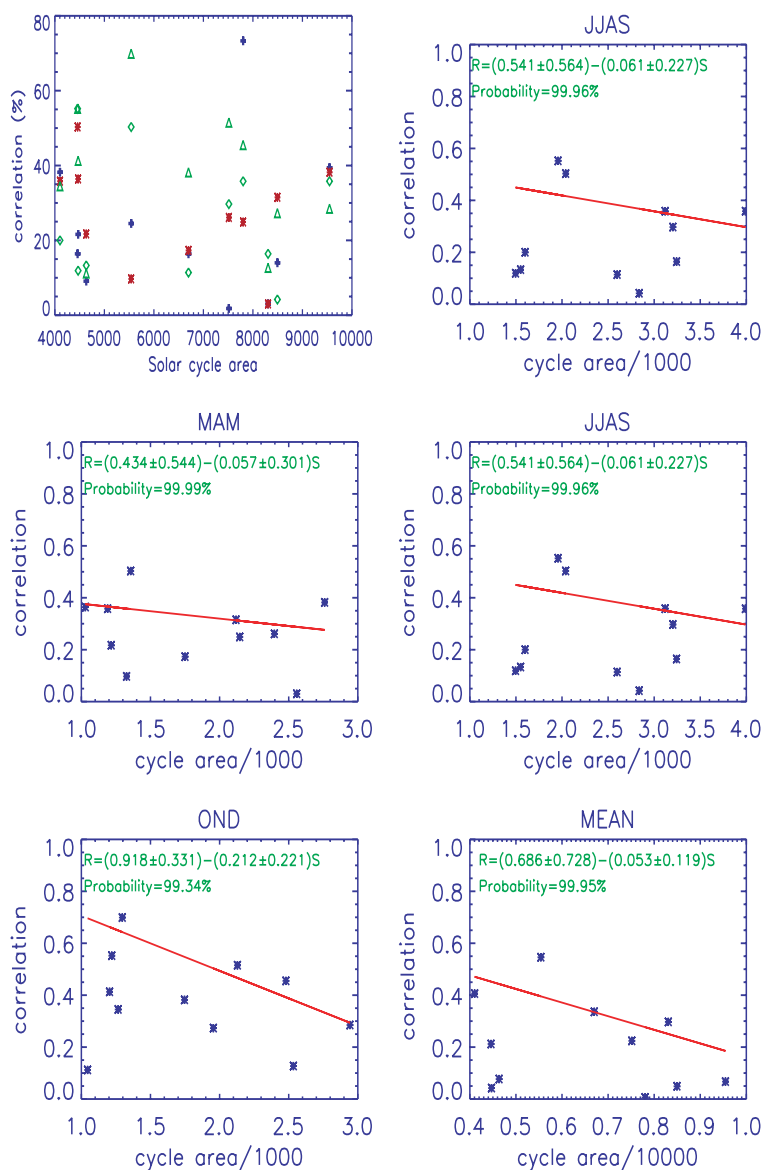


Figure 1. Different cycle-to-cycle variations of the correlation coefficients for the seasonal months are plotted against the annual area of the solar cycle. Different symbols in the first plot are: JF as +, MAM as *, JIAS as $\diamond$ and OND as $\triangle$. The rest of the plots represent the same as in the first plot, except that data are presented separately (for different seasonal months) with a line (in red colour) that is obtained from the linear least square fit is also over plotted.

overall trend that irrespective of signs of the correlation coefficients, the monsoon rainfall variability is strongly correlated during the weak sunspot activity and weakly correlated during the strong sunspot activity. That means during the 11-year solar cycle, *a weak sunspot activity leads to a strong Indian rainfall variability and vice versa.*

3. Conclusions and discussion

From the correlative analysis of 130 years data of the sunspot and the Indian monsoon rainfall activities, conclusions of the present study are:

- with a moderate-to-high significance sunspot activity is correlated with the Indian rainfall activity,
- irrespective of signs of the correlation coefficients, for a particular period of a solar cycle, there appears to be an overall trend that rainfall variability is high during the low solar activity than the high solar activity.

As for the low-to-moderate correlations, as we discussed in the previous study (Hiremath and Mandi 2004), there is a *causal relationship* between the sunspot activity, galactic cosmic rays (GCR) and the rainfall in the following way. The GCR activity is the source of ions in the Earth's atmosphere. We know that the condensation of water vapour into water drops is mediated by the ions in the atmosphere. Thus any change in the GCR activity correspondingly affects the rainfall variability. As the intensity of the GCR is inversely proportional to the solar activity, increase in solar activity results in reducing the intensity of the GCR flux. This ultimately results in both reducing the activity in nucleation of the cloud particles and suppressing the rainfall variability (Parker 1999). That means during high sunspot activity we should get a small correlation coefficient (considered as absolute value) between the occurrence of sunspot and the rainfall variabilities. That is what we are finding in most of the plots.

It is expected from the previous study (Hiremath and Mandi 2004) that for all the seasonal and annual rainfall data one should get positive correlation. Whereas negative correlations in some of the cycle periods could be due to the changing spatial and the temporal occurrence pattern of the aerosol particles either due to intermittent volcanic eruptions (Robock 2000) or due to intrusion of interstellar dust particles in the atmosphere. The recent study (Vinoj *et al.* 2004) shows the rain-deficit of the summer monsoon due to large aerosol optical depths. In fact there is a constant accretion ($\sim (4.0 \pm 2.0)10^6$ kg/year) of cosmic dust due to extra terrestrial origin on the Earth's atmosphere (Lone & Brownlee 1995; Yada *et al.* 2000; Lal & Jull 2002). In the present context, it is very difficult to judge as to which of the aerosol particles influences the Indian rainfall.

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Gravitational Clustering of Galaxies in an Expanding Universe

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Abstract. We inquire the phenomena of clustering of galaxies in an expanding universe from a theoretical point of view on the basis of thermodynamics and correlation functions. The partial differential equation is developed both for the point mass and extended mass structures of a two-point correlation function by using thermodynamic equations in combination with the equation of state taking gravitational interaction between particles into consideration. The unique solution physically satisfies a set of boundary conditions for correlated systems and provides a new insight into the gravitational clustering problem.

Key words. Gravitational clustering—thermodynamics—correlation function: structure of universe.

1. Introduction

Galaxies interact gravitationally and the characterization of this clustering is a problem of current interest. The gravitational interaction between galaxies and galaxy clusters have played an important role in the evolution of the observed universe. Various theories of the cosmological many body problem have been developed mainly from a thermodynamic point of view. We have made use of the equations of state along with the correlation functions for the development of a theoretical model. This can be done by solving a system of Liouville's equation or BBGKY-hierarchy equations and have been discussed by workers like Saslaw (1972) and Peebles (1980). But BBGKY hierarchy equations are too complicated to handle for higher order correlation functions. However, the lowest order two-point correlation function is useful for discussing the phenomenon of gravitational clustering of galaxy clusters which contain information on all the higher n -particle correlations in the full BBGKY-hierarchy (Peebles 1980; Zhan & Dyer 1989; Hamilton 1993). The physical validity of the application of thermodynamics in the clustering of galaxies and galaxy clusters has been discussed on the basis of N -body computer simulation results (Itoh *et al.* 1993). In gravitational thermodynamics the value of b which is the ratio of gravitational correlation energy to twice kinetic energy measures two-point correlation function ξ_2 and depends on the average number density $\bar{n}$, temperature T and the interparticle distance r . Thus it is valuable to understand the functional form of ξ_2 which depends upon the value of $b(n, T)$. The

gravitational galaxy clustering carried out by Ahmad *et al.* (2002, 2006) ensures a more fundamental statistical mechanical description of the cosmological many-body problem.

The two partial differential equations developed for the galaxy clusters (with point mass and extended mass structures) in an expanding universe provide a new approach for understanding the phenomena.

2. Approaches

2.1 Development of differential equation for two-point correlation function

Taking the lead of Saslaw & Hamilton (1984), we start with a general pair of equations of state for the internal energy U and the pressure P :

$$U = \frac{3}{2}NT(1 - 2b), \quad (1)$$

$$P = \frac{NT}{V}(1 - b). \quad (2)$$

These two equations represent the equations of state (e.g., Hill 1956), where the dimensionless

$$b = \frac{2\pi Gm^2\bar{n}}{3T} \int_V \xi_2(\bar{n}, T, r) r dr \quad (3)$$

defines the measuring correlation parameter. Here $\bar{n} = N/V$ is the average number density of the system of particles each of mass m . T is the temperature, V the volume, G is the universal constant of gravitation, $\xi_2(\bar{n}, T, r)$ is the two-particle correlation function and r the distance. These expressions assume a large volume V for their validity.

The thermodynamic parameters P , V and T are related by

$$P = P(V, T) \quad \text{and} \quad U = U(V, T).$$

According to the first law of thermodynamics, for any system

$$dQ = dU + PdV \quad (4)$$

where dQ is the amount of heat flowing into the system, dU is the gain in its internal energy and PdV is the work done by the system as its volume increases by the infinitesimal amount dV under the internal pressure P .

For fixed number of particles N , the internal energy U is related to V and T as

$$dU = \left(\frac{\partial U}{\partial V} \right)_T dV + \left(\frac{\partial U}{\partial T} \right)_V dT. \quad (5)$$

Hence equation (4) becomes:

$$dQ = \left[\frac{\partial U}{\partial V} + P(V, T) \right] dV + \left(\frac{\partial U}{\partial T} \right)_V dT. \quad (6)$$

According to the second law of thermodynamics, if a system undergoes an infinitesimal quasi-static change as a result of which it gains an infinitesimal amount of heat dQ at the absolute temperature T , then there exists a perfect differential dS of a function S (entropy) defined by:

$$dS = \frac{dQ}{T}. \quad (7)$$

It must be understood that this definition applies only if the system undergoes an infinitesimal quasi-stationary change.

The combination of equation (7) with equation (6) leads to

$$dS = \frac{dQ}{T} = \frac{1}{T}(dU + PdV) = \frac{1}{T} \left[\frac{\partial U}{\partial V} + P(V, T) \right] dV + \left(\frac{\partial U}{\partial T} \right) \frac{1}{T} dT. \quad (8)$$

As dS is a perfect differential, we must have

$$\frac{\partial}{\partial T} \left[\frac{1}{T} \left(\frac{\partial U}{\partial V} \right) + \frac{P}{T} \right] = \frac{\partial}{\partial V} \left[\left(\frac{\partial U}{\partial T} \right) \frac{1}{T} \right] \quad (9)$$

which leads to:

$$\left(\frac{\partial U}{\partial V} \right)_{T,N} = T \left(\frac{\partial P}{\partial T} \right)_{V,N} - P. \quad (10)$$

The probability for finding a galaxy in volume dV_1 and also one in volume dV_2 is related to two-point correlation function $\xi_2(r)$ by:

$$P_{12} = \bar{n}^2 [1 + \xi_2(r)] dV_1 dV_2. \quad (11)$$

In general ξ_2 will depend on the absolute positions $\mathbf{r}_1$ and $\mathbf{r}_2$ of two volume elements. However, if we average overall positions and all directions, then ξ_2 will become a function only of the separation $\mathbf{r} = |\mathbf{r}_1 - \mathbf{r}_2|$. For $\xi_2(r)$ to provide a good description, the system must be statistically homogeneous so that on an average overall, the galaxies in a given volume $V (= \frac{4}{3}\pi r^3)$ will be the same as for any sufficiently large subset of galaxies in the same volume. In the grand canonical ensemble, the two-point correlation function ξ_2 depends on $\bar{n}$, T as well as on r . Thus we may write

$$\xi_2 = \xi_2(\bar{n}, T, r), \quad (12)$$

$$d\xi_2 = \frac{\partial \xi_2}{\partial \bar{n}} d\bar{n} + \frac{\partial \xi_2}{\partial T} dT + \frac{\partial \xi_2}{\partial r} dr, \quad (13)$$

$$\frac{d\xi_2}{dV} = \frac{\partial \xi_2}{\partial \bar{n}} \frac{d\bar{n}}{dV} + \frac{\partial \xi_2}{\partial T} \frac{dT}{dV} + \frac{\partial \xi_2}{\partial r} \frac{dr}{dV}. \quad (14)$$

Using the condition

$$\frac{dT}{dV} = 0. \quad (15)$$

Using equations (1) to (3) in equation (10), we have:

$$3\bar{n} \frac{\partial \xi_2}{\partial \bar{n}} + T \frac{\partial \xi_2}{\partial T} - r \frac{\partial \xi_2}{\partial r} = 0. \quad (16)$$

Equation (16) is a first order partial differential equation for two-point correlation function. It is characterized by number density $\bar{n}$, temperature T and the interparticle distance r , therefore two-point correlation function ξ_2 will depend on the values of $\bar{n}$ and T for the ensemble as well as on the spatial co-ordinate r in a statistically homogeneous distribution of galaxies clustering gravitationally in an expanding universe.

The galaxies in an expanding universe are treated as point mass particles, but actually they have extended structures and therefore make use of a softening parameter ϵ .

Hence for extended mass distribution, the internal energy U and pressure P can be written as,

$$U_{\text{ext}} = \frac{3}{2}NT - \frac{2\pi Gm^2 N^2}{3V^2} \int_V \xi_2(\bar{n}, T, r) \left[1 + \frac{\epsilon^2}{r^2} \right]^{\frac{-1}{2}} \frac{dV}{4\pi r}, \quad (17)$$

$$P_{\text{ext}} = \frac{NT}{V} - \frac{2\pi Gm^2 N^2}{3V^2} \int_V \xi_2(\bar{n}, T, r) \left[1 + \frac{\epsilon^2}{r^2} \right]^{\frac{-3}{2}} \frac{dV}{4\pi r}. \quad (18)$$

Again repeating the same procedure the new derived equation is,

$$3\bar{n} \frac{\partial \xi_2}{\partial \bar{n}} + T \frac{r^2}{(r^2 + \epsilon^2)} \frac{\partial \xi_2}{\partial T} - r \frac{\partial \xi_2}{\partial r} = 0. \quad (19)$$

Equation (19) is a first order partial differential equation for a two-point correlation function of galaxies with extended structures clustering gravitationally. For extended mass structures, the value of ϵ is taken between 0.01 and 0.05 (in the units of total radius). The thermodynamic description of a two-point correlation function for describing the gravitational interaction of galaxies can be defined by the physical behaviour of equations (16) and (19) respectively.

2.2 Functional form of two-point correlation function

The partial differential equations developed in (16) and (19) can have a variety of solutions, but we are interested in a solution that is physically valid. This can be verified when a set of boundary conditions assigned for two-point correlation function are satisfied.

- The gravitational clustering of galaxies in a homogeneous universe requires ξ_2 to have a positive value which obviously depends upon the limiting values of $\bar{n}$, T and r .
- When $\bar{n}$, T and r are very small (approximately tending to zero), the two-particle correlation function ξ_2 will increase except for the number density $\bar{n}$. Similarly when $\bar{n}$, T and r are very large, the corresponding correlation function ξ_2 will decrease except for the number density $\bar{n}$.
- When two-particle correlation function increases, the clustering of galaxies becomes dominant because of virial equilibrium, which suggests that at low

temperatures and high densities more and more clusters are formed. Thus we can write this boundary condition as:

When $\bar{n}T^{-3}$ is very large (approximately tending to infinity), the two-particle correlation function ξ_2 will increase and the measuring correlation parameter b approaches to 1. The reverse will happen for $\bar{n}T^{-3}$ to be very small (approximately tending to zero).

Equation (16) can be integrated along with its characteristics as,

$$\frac{d\bar{n}}{3\bar{n}} = \frac{dT}{T} = \frac{dr}{-r}. \quad (20)$$

This equation can have a variety of solutions in different combinations like:

$$\xi_2 = f(\bar{n}T^{-3}, Tr) \quad \text{or} \quad f(Tr, \bar{n}r^3) \quad \text{or} \quad f(\bar{n}r^3, \bar{n}T^{-3}).$$

All these functions are scale invariant under the transformation

$$\bar{n} \rightarrow \lambda^{-3}\bar{n}, \quad T \rightarrow \lambda^{-1}T \quad \text{and} \quad r \rightarrow \lambda r$$

where λ is a constant.

We are interested in a solution which satisfies the above-mentioned boundary conditions.

After trying many combinations, we choose the solution as

$$\xi_2(\bar{n}, T, r) = \left[\frac{C_1 \bar{n} T^{-3}}{1 + C_1 \bar{n} T^{-3}} \right]^2 \frac{1}{C_1 \bar{n} T^{-3}} \frac{1}{(C_2 Tr)^2} \quad (21)$$

and

$$\xi_2(\bar{n}, T, r) = \frac{b(1-b)}{C(Tr)^2} \quad (22)$$

C , C_1 and C_2 are constants. On testifying the boundary conditions, we see that the two-particle correlation function defines its exact behaviour. The unique solutions shown in equations (21) and (22) show the dependence of ξ_2 on $\bar{n}T^{-3}$, Tr and b . When we consider other combinations shown in the ξ_2 solutions, two-particle correlation function again shows its dependence on $\bar{n}$, T and r . The unique solution of equation (16) can also be verified by using the value of the parameter b defined by Saslaw & Hamilton (1984) which is described by equation (3). Rewriting equation (3) as,

$$b = \frac{2\pi Gm^2 \bar{n}}{3T} \int_V \xi_2(\bar{n}, T, r) \frac{dV}{4\pi r}. \quad (23)$$

On substituting equation (21) in (23), the result leads to

$$b = \frac{2\pi Gm^2}{9C_1 C_2^2} \frac{C_1 \bar{n} T^{-3}}{1 + C_1 \bar{n} T^{-3}} = \frac{\beta \bar{n} T^{-3}}{1 + \beta \bar{n} T^{-3}}. \quad (24)$$

This is the same equation as used earlier by Saslaw & Hamilton (1984) for the thermodynamic dependence of b . The self evaluation of b clarifies some important consequences:

- The unique solution chosen in equation (21) of a two-particle correlation function gets immediately confirmed.
- The thermodynamic parameter b shows a complete dependence on the combination $\bar{n}T^{-3}$.

2.3 Power law for ξ_2

The two-point correlation function in a gravitational galaxy clustering obeys power law Peebles (1980), which has also been confirmed from N -body computer simulations (Itoh *et al.* 1993) and is written as

$$\xi_2 = r^{-1.8}. \quad (25)$$

From equation (3), we can write

$$\frac{\partial \bar{n}}{\partial V} \left[\frac{\partial b}{\partial \bar{n}} - \frac{b}{\bar{n}} \right] = \frac{2\pi G m^2 \bar{n}}{3T} \frac{\xi_2(\bar{n}, T, r)}{4\pi r}. \quad (26)$$

The combination of physical limits and mathematical simplicity made by Saslaw & Hamilton (1984) suggests that the simple form of b is described by equation (24). β is a positive dimensionless constant and is independent of $\bar{n}$, T and r .

$$\frac{\partial b}{\partial \bar{n}} = \frac{b(1-b)}{\bar{n}}, \quad \frac{\partial \bar{n}}{\partial V} = \frac{-\bar{n}}{V}. \quad (27)$$

The substitution of equation (27) in (26) gives the modified form of equation (25) as:

$$\xi_2 = r^{-2.0} \quad (28)$$

which shows that the power law for a two-point correlation function developed is in close agreement with Peebles law, and hence confirms the applicability of gravitational quasi-equilibrium thermodynamics used to study the two-point correlation function for understanding gravitational clustering of galaxies in an expanding universe.

3. Discussion

Many theories have been discussed so far for understanding the gravitational clustering of galaxies in an expanding universe, but in our approach we have made use of basic thermodynamics, correlation functions and have found their applicability in clarifying many important results. The development of equations (16) and (19) indicates the dependence of ξ_2 on $\bar{n}$, T and r . The results of two-point correlation function ξ_2 can be extended to higher order correlations also. The functional form of ξ_2 has been obtained from various physical and boundary conditions. Our results also indicate the dependence of b on the specific combination $\bar{n}T^{-3}$ and therefore clarifies the earlier result of Saslaw & Hamilton (1984). Till date, all the theories on gravitational interaction of correlated systems (e.g., galaxies, clusters of galaxies, etc.) are confined to point mass systems only. However, the theory so developed is not only applicable to point mass systems but can also be applicable to extended structures, thus indicating the dependence on one more parameter called softening parameter ϵ on ξ_2 . A theoretical analysis of a thermodynamic model along with the correlation functions used will lead to a new direction for understanding a large scale structure of the universe.

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QSO Pairs *across* Active Galaxies: Evidence of Blueshifts?

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Abstract. Several QSO pairs have been reported and their redshifts determined, where the two objects in each pair are located *across* an active galaxy. The usually accepted explanation of such occurrences is that the pair is ejected from the parent galaxy. Currently interpreted redshifted spectra for both the QSOs imply that both the objects are receding from the observer. However, ejection can occur towards and away from the observer with equal probability. We argue that for a system with two QSOs lying *across* the parent galaxy, ejection should have occurred in opposite directions, whereby one object will be approaching us and the other will be receding from us. The former would exhibit a blueshifted spectrum. We analyse here a sample of four such pairs and show that the observed spectrum of one QSO in each pair can be interpreted as blueshifted. The other exhibits the usual redshifted spectrum. A scenario based on the ‘sling-shot’ mechanism of ejection is presented to explain the occurrences of the pairs in opposite sides of the active galaxies moving in opposite directions.

Key words. Quasars: emission lines—quasars: absorption lines—cosmology: miscellaneous.

1. Introduction

A literature search would reveal that there is a growing number of Quasi Stellar Object (QSO) pairs being observed, where the two objects in the pair are located *across* an active galaxy. We present here a sample of four such pairs and their associated active galaxies (Table 1). In Table 1, column (1) gives the name of the active galaxy and the reference, column (2) gives its redshift, columns (3) and (5) give the names of the two QSOs forming the pair across the active galaxy, and columns (4) and (6) give the redshifts of the two QSOs respectively. In each of these pairs, two QSOs have been identified and their redshifts determined, which are much larger than the redshift of the parent galaxy. Ejection from the parent galaxy is the explanation of such occurrences.

However, the basic mechanism of ejection necessitates that the process can occur in all directions with equal probability for randomly ejected objects. Objects being ejected may therefore move away from us as well as towards us, again, with equal probability. Such a situation is specially important and significant when the two objects are located *across* the parent galaxy, implying that the objects have been ejected by the galaxy to its two opposite sides, and hence, in two *opposite directions*. An ejection

Table 1. QSOs across galaxies and redshifts.

Galaxy (1)	z_r (2)	QSO (3)	z_r (4)	QSO (5)	z_r (6)
NGC 3628(i)	0.0028	1WGAI 1120.2 + 1332	0.995	1WGAI 1120.4 + 1340	0.981
NGC 3842(ii)	0.0260	1141.8 + 2013 (QSO 1)	0.335	1141.4 + 2014 (QSO 2)	0.946
NGC 5548(iii)	0.0170	1E 1415.1 + 2527	0.560	1E 1416.7 + 2526	0.674
Mark 205(iv)	0.0700	1218.7 + 7522	0.645	1219 + 7533	0.460

References: (i) Arp *et al.* (2002); (ii) Arp (1984); (iii) Arp (1997); (iv) Arp (1995).

towards the observer would, of course, exhibit blueshifts in their observed spectra, in at least some cases under suitable conditions. In this respect, the kinematics of the ejection mechanism suggests that at *larger* redshifts (≥ 3.5), “randomly directed motions will produce almost an order of magnitude more redshifts than blueshifts”, and “as the mean speed of the sources increases, the fraction of blueshifts decreases” (Gordon 1980). This implies that for objects with *smaller* redshifts moving at slower speeds, the fraction of blueshifts to redshifts is higher than that for high redshift objects moving at larger speeds. All QSO pairs in our sample have redshifts smaller than 1.0, much smaller than the redshift of “ ≥ 3.5 ” (Gordon 1980), and, hence, the probability of some of them being blueshifted is quite high. Furthermore, as commented by Popowski & Weinzeirl (2004), “it is unrealistic to expect that the ejection process would always meet the conditions to produce only redshifts”. As such, at least a certain fraction of the vast population of QSOs now known is expected to exhibit blueshifts.

Nevertheless, the spectra of both objects in each pair have been interpreted as redshifted. Ejection has thus been considered away from us for *both* objects, without paying due consideration to the probability of the ejection towards the observer, apparently with the assumption that blueshifts are not observable in QSO spectra. The reality, however, is that no attempt is made in interpreting observed spectra in terms of blueshifts. The line identification program is such oriented as to ensure that all spectra are, almost as a rule, interpreted as redshifted.

We suggest that the two objects in each QSO pair under investigation here have been ejected from the parent active galaxy in opposite directions, resulting in one approaching us and the other receding from us, which would result in one of the objects exhibiting a blueshifted spectrum. The purpose of the present paper is to show that the spectrum of one object in each pair of the QSOs can indeed be interpreted as blueshifted.

2. Blueshifted spectra

Putsil’nik (1979) demonstrated that the most frequently used UV lines CIII] 1909 and MgII 2798 for redshift identification of observed emission lines in QSOs may actually be misidentifications, and the IR lines of the Paschen series, *viz.*, $P\beta$ 12818 and $P\alpha$ 18751 respectively, may be better identifications instead, producing ‘violet shifts’. Additionally, a literature search reveals many inconsistencies and inadequacies in redshift identifications. Thus, expected strong features are sometimes not seen at expected positions or observed as very weak lines or of unusual profiles, the redshift computed on the basis of observed lines may not be able to interpret additional features

seen in subsequent observations, problems appear in energetics based on the calculated redshift values. We are, therefore, concerned about the correctness of the identification process, and, hence, about the possibility of misidentification of observed lines in the redshift determination of QSOs and other extragalactic objects. This necessitates alternative identifications and may lead to blueshifts.

It has been shown in recent years that blueshifts can interpret observed emission lines of QSOs, and, from various considerations, such interpretations in many cases are more convincing than the redshift interpretations (Basu & Haque-Copilah 2001). Very unusual spectra of three additional QSOs, *viz.*, SDSS 1533–00, PG 1407 + 265 and PKS 0635–752, emission and absorption, which cannot be explained properly by the redshift mechanism, have been interpreted in terms of blueshifts (Basu 2004). Spectra of two active galactic nuclei (AGN), *viz.*, PKS 2149–306 and CXOCDFS J033225.3–274219, each exhibiting an X-ray emission feature, which cannot be explained by the redshift determined from its optical spectrum, have been successfully interpreted in terms of blueshifts (Basu 2006a). Blueshifts have also been demonstrated to explain observed spectra of other extragalactic objects, *viz.*, several high redshift galaxies (Basu 1998), the galaxy STIS123627 + 621755 (the redshift interpretation being unable to explain this spectrum) (Basu 2001a), host galaxies of several supernovae Ia (Basu 2000), host galaxies of several gamma ray bursts (Basu 2001b), and also the puzzling spectra of the galactic X-ray source 1E 1207.4–5209 associated with the SNR G296.5 + 10.0 (Basu 2006b). In all the above cases, serious inconsistencies exist in redshift interpretations of the spectra of the objects concerned. It is therefore possible that redshifts have been assigned to spectra of some extragalactic objects due to misidentification of observed lines.

3. Results

We re-examined the observed spectra of the eight QSOs of the four pairs in our sample, and found several examples of incompleteness and inadequacy in the redshift identification process of some of the objects, as discussed in section 4. As such, we have interpreted the spectra of one object in each pair as blueshifted by identifying the observed lines with search lines of longer wavelengths, and have determined the blueshift values of each line and, hence, of each such object. We have followed the standard procedure of the identification process for the blueshift determination, *viz.*, at least two observed lines must exhibit the same ‘shift’ (red or blue), whether emission or absorption, when identified with two separate search lines, and the same ‘shift’ must be exhibited for any third or more lines (Basu 1973a, 1973b). Furthermore, when necessary, the higher order line(s) of a series (Balmer or Paschen) has been identified, the lower order line(s) being outside the observed region of the spectrum. Also, in some cases, lower order line(s) of a series has been identified, the higher order line(s) may be too weak to be seen. Results are shown in Table 2.

In Table 2, column (1) is the name of the active galaxy, column (2) gives the name of the QSO, whether the features are emission (EM) or absorption (ABS) and the reference of its spectrum, column (3) is the observed wavelength of the line in the spectrum (λ_o), column (4) is the wavelength of the search line identified with the observed line for determining the redshift (λ_r), column (5) is the redshift value of the line (z_r), column (6) is spread in redshift (Δz_r) measured as the difference between the maximum and minimum values of redshifts in the system, column (7) is the wavelength of the search

Table 2. Redshifts and blueshifts in observed lines of five QSOs.

Galaxy (1)	QSO (2)	λ_o (3)	λ_r (4)	z_r (5)	Δz_r (6)	λ_b (7)	z_b (8)	Δz_b (9)
NGC 3628	1120 + 13 EM (i)	3720	CIII] 1909	0.9800	0.008	P β 12818	0.7098	0.0060
		5565	MgII 2798	0.9880		P α 18751	0.7032	
NGC 5548	1416 + 25 EM (ii)	4684	MgII 2798	0.6740		P α 18751	0.7502	
Mark 205	1218 + 75 EM (iii)	4600	MgII 2798	0.6440	0.012	H β 4861	0.0537	0.0060
		6173	[OII] 3727	0.6564		H α 6563	0.0594	
	ABS (iii)	4275	FeII 2587	0.6525	0.011	H2 20338	0.7898	0.0087
		4307	FeII 2600	0.6565		H2 20587	0.7908	
		4600	MgII 2796	0.6452		H2 21218	0.7832	
		4627	MgII 2803	0.6507		H2 22233	0.7919	
NGC 3842	1141 + 20 (QSO 2) EM (iv)	3714	CIII] 1909	0.9460	0.003	H β 4861	0.2360	0.0100
		4050	?	?		[CaV] 5301	0.2370	
		4494	?	?		HeI 5876	0.2350	
		4804	[OII] 2470	0.9450		[OI] 6300	0.2375	
		5444	MgII 2798	0.9457		HeI 7065	0.2294	
		8455	H γ 4340	0.9482		P γ 10938	0.2270	
	ABS (iv)	3528	SiII 1808–1817	0.9510	0.023 0.019	SiII 5454	0.3531	0.0183
		4300?	?	?		H α 6563	0.3348	
		8369?	H γ 4340?	0.9280		P β 12818	0.3471	

References: (i) Arp *et al.* (2002); (ii) Stocke *et al.* (1983); (iii) Gioia *et al.* (1984); (iv) Arp (1984) (see also Basu & Haque-Copilah 2001).

line identified with the observed line for determining the blueshift (λ_b), column (8) is the blueshift value of the observed line (z_b), column (9) is the spread in blueshift (Δz_b) measured as the difference between the maximum and minimum values of blueshifts in the system.

The mean blueshifts of the four objects, from Table 2, are 0.7065, 0.7502, 0.0566 and 0.2340, respectively from top downwards. Additionally, two of the QSOs, *viz.*, 1218 + 75 and 1141 + 20 (QSO2), exhibit absorption lines as well, and these have been identified with search lines of longer wavelengths to obtain mean blueshifts 0.7890 and 0.3450 for the two objects respectively.

Absorption lines are known to be formed partly by hydrogen clouds at various stages of evolution to galaxy formation and partly by metals produced in haloes of galaxies.

4. Discussion

We have identified the two emission lines at 3720 Å and 5565 Å in the QSO 1120 + 13 with the IR lines P β 12818 and P α 18751, instead of the UV lines CIII] 1909 and MgII 2798 respectively. The possibility that these two IR lines are better identifications of some observed lines rather than the UV lines was suggested by Putsil'nik (1979), as mentioned earlier (section 2).

The QSO 1416 + 25 exhibits only one emission line, *viz.*, that at 4684 Å, which has been identified with MgII 2798 in the redshift interpretation. In principle, it can be identified with any search line, and the proper identification needs at least two lines, as explained above. However, in the absence of any other line and any other information of this line, and following Putsil'nik (1979), we have identified the line with P α 18751. The other possible alternative is, of course, H α 6563, but that would require a much stronger line. On the other hand, H β 4861 is ruled out in this case as that would bring H α 6563 within the observed region which is not seen.

For the QSO 1218 + 75, the blueshift identification of the line at 4600 Å with H β 4861 is justified as the H α 6563 expected within the observing region is seen at 6173 Å. Both are fairly strong lines. The absorption lines are identified in the blueshift interpretation with four of the strongest molecular hydrogen (H₂) lines.

The emission lines of the QSO 1141 + 20 (QSO 2) have been shown earlier to be better interpreted as blueshifted (Basu & Haque-Copilah 2001). But the spectrum exhibits absorption lines also, and here we treat the complete spectrum – both emission and absorption. On the other hand, the redshift interpretation of the spectrum of 1141 + 20 (QSO 2) is very confusing. It has been described as “difficult to identify” for the redshift measurement, with “unusual features”, and “H γ and H α in emission, if present at all, are very weak”. Two emission lines at 4050 Å and 4494 Å, and the broad “so strong” absorption line at 4300 Å can not be identified in the redshift interpretation. In addition, the absorption line at 3528 Å has been identified with a multiple of lines *viz.*, SiII 1808–1817. We have identified *all* the observed lines in this object in the blueshift interpretation – the above two emission lines as [CaV] 5301 and HeI 5876 respectively, and the broad absorption line as the H α 6563, the strongest of the search lines.

It should be noted here that the blueshift values presented in this paper are results of superposition of the cosmological redshifts and the Doppler shifts due to the object approaching the observer.

5. The ‘spread’ statistic: Goodness of fit

The quantity spread, Δz in Table 2 (columns 6 and 9), is a measure of the goodness of fit for the identification of the observed lines with search lines for computing the ‘shift’, red or blue, for a system. In principle, Δz should be as close to zero as possible. However, in practice, the quantity depends on, and is very sensitive to, the value of the observed wavelength (λ_o , column 3, Table 2) which is used to evaluate the ‘shift’. Unfortunately, the exact determination of λ_o is very difficult in practice, even in high s/n and high resolution spectra, as the line profile may be broad, double- or multi-peaked, with complicated structure, blending, etc. Physically, these may involve net flows, screening, gradients, partial absorption, etc. Local velocity drifts, infall or outflow of materials (yet undecided, Penston *et al.* 1990), etc. may also move λ_o away from the expected position. Uncertainties in the determination of λ_o due to the above reasons may, in some cases, result in somewhat larger values of Δz , in both redshifts and blueshifts. At least up to a certain extent, this should not be considered as errors in measurement, but may have physical reasons. Having said that, Table 1 would show that, in general, Δz_b , i.e., spreads in blueshift values, are somewhat smaller, and hence closer to zero, than Δz_r , i.e., spreads in redshift values.

6. A possible scenario: Ejection mechanism

The analysis presented here shows that QSO pairs *across* active galaxies comprise two objects – one approaching us and thus exhibiting blueshifts, the other moving away from us and thus exhibiting redshifts. Such a scenario can be achieved in terms of the ejection mechanism. Merger of black holes is known to lead to their ejections in opposite directions by the so-called ‘sling-shot’ mechanism when the system becomes unstable (Saslaw *et al.* 1974; Valtonen 1976a, 1976b). Merging of galaxies, each hosting a supermassive black hole at its centre, which are known to be seats of activities (Basu *et al.* 1993; Capetti *et al.* 2005), may produce such a situation. Initially, a binary system is formed by the two central black holes (Valtoja *et al.* 1989), and such a system has been reported to be detected in NGC 6240 (Komossa *et al.* 2003). As the merger process proceeds further, the supermassive black holes (primaries) are ejected at relativistic or non-relativistic speeds “in two opposite directions” (Mikkola & Valtonen 1990). Evidence of ejection of a supermassive black hole by the ‘sling shot’ mechanism resulting from merger of galaxies has recently been presented (Haehnelt *et al.* 2006). Additionally, satellite black holes of intermediate masses are also believed to accompany the central supermassive black holes in galaxies (Carr 1978; Carr *et al.* 1984) and are also ejected, some of them assuming eccentric orbits around the primaries (Valtonen & Basu 1991).

It is further known that a black hole at the centre of a galaxy possesses a gaseous accretion disk around it, and this is believed to survive the tidal disruption accompanying the ejection process (Rees & Saslaw 1975; Lin & Saslaw 1977; De Young 1977). A QSO may be formed by the interaction between the disk with the black hole and the surrounding (Rees 1984; Osterbrock & Mathews 1986; Valtonen & Basu 1991; Spriegel 2005). One can further envisage that the same process of interaction would occur between the surroundings and the gaseous disks around the satellite black holes, as in the case of their primary counterparts, although at reduced scales as their masses

are smaller, and eventually some of the satellite black holes may end up as faint or nascent or other galaxies.

The merger of two galaxies thus finally gives birth to two QSOs ejected in opposite directions. Each QSO may be accompanied by one or more galaxy-like objects, and the latter act as absorbing clouds when aligned along the line of sight. It may be noted in this connection that absorbing clouds have been shown earlier to be linked with the birth of QSOs themselves (Basu 1982). Moreover, faint, nascent and other galaxies associated with QSO-like objects have been reported to be observed (Dressler *et al.* 1993; Tripp *et al.* 1998; Tresse *et al.* 1999).

It can therefore be conceived that QSO pairs *across* active galaxies are the results of the ejection process, *viz.*, the ‘sling shot’ mechanism. One of the pairs is approaching us exhibiting blueshifts in emission and absorption lines, the latter, when observed, being produced by the accompanying absorbing clouds in the form of faint or nascent or other galaxies ejected by the same process but at larger speeds and hence showing larger absorption blueshifts than the corresponding emission blueshifts. The other is receding from us exhibiting the usual redshifted spectra.

7. Concluding remarks

We have followed an entirely different path, *viz.*, blueshifts in extragalactic spectra, and not a variant of existing ones, to interpret observational data. We are motivated by the possibility that blueshifts have been ignored by the astronomical community in the modern line identification programs. This, in its turn, has led to wrong approaches. One such example is the assumption that the ejection mechanism would always occur away from the observer, thus obeying all the rules of redshifts only. We have shown here that blueshifts are real possibilities, and observations obey the basic notion of the ejection process, *viz.*, objects should be ejected in all directions with equal probability.

Finally, it should be emphasized that blueshifts and redshifts are complementary and not contradictory. Modern observations would bring in many unusual systems some of which may not be explained by the usual redshift hypothesis, and blueshifts may explain them. Ignoring blueshifts may result in missing some important cosmological scenarios.

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Report on the Dynamical Evolution of an Axially Symmetric Quasar Model

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Abstract. The role of the angular momentum in the regular or chaotic character of motion in an axially symmetric quasar model is examined. It is found that, for a given value of the critical angular momentum L_{zc} , there are two values of the mass of the nucleus M_n for which transition from regular to chaotic motion occurs. The $[L_{zc}-M_n]$ relationship shows a linear dependence for the time independent model and an exponential dependence for the evolving model. Both cases are explained using theoretical arguments together with some numerical evidence. The evolution of the orbits is studied, as mass is transported from the disk to the nucleus. The results are compared with the outcomes derived for galactic models with massive nuclei.

Key words. Galaxies: orbits—regular and chaotic motion.

1. Introduction

High resolution data obtained by the Hubble Space Telescope (HST) indicate that nuclear activity may be a common phenomenon during the lifetime of a galaxy. Furthermore, it is generally believed that the accretion episodes and the emitted nuclear power strongly depend on the mass of the system. One of the unsolved problems in Astronomy is the problem of the energy of quasars. Most astronomers believe that quasars are supermassive stellar objects (SMOs) with huge luminosity. This immense luminosity is the result of the accretion of matter outside those SMOs (see e.g., Cavaliere & Padovani 1988; Dobrzycki & Bechtold 1991; Siemiginowska 1991; Canizzo 1993; Peng & Chou 1997, 2001).

In the present paper we use a simple dynamical model for a quasar, in order to study the properties of motion. We believe that this is of interest because such studies are not very common in galactic dynamics. This study will give useful information for the dynamical properties and evolution of quasars. On the other hand, we will be able to compare the dynamical behavior of these systems with the results earlier obtained for galaxies and non-axially symmetric quasar models (see Papadopoulos & Caranicolas 2005).

We are particularly interested to find the nature of motion (regular or chaotic) and to seek for relationships connecting chaos and important physical parameters such as

mass and angular momentum. As this was already done for galaxies, we will be able to compare the present results with the already well known results for galaxies (see Caranicolas & Innanen 1991; Caranicolas & Papadopoulos 2003).

Our dynamical model has two components. A disk and a massive nucleus and it is given by the following equation:

$$V(r, z) = -\frac{M_d}{[(\alpha + (z^2 + h^2)^{1/2})^2 + r^2]^{1/2}} - \frac{M_n}{(r^2 + z^2 + c_n^2)^{1/2}}, \quad (1)$$

where r, z are cylindrical co-ordinates, M_d is the mass, α is the scale length of the disk and h corresponds to the disk scale height. M_n is the mass, while c_n represents the scale length of the nucleus. Here again, the well known system of galactic units is used, where the unit of length is 1 kpc, the unit of time is 0.97746×10^8 yr and the unit of mass is $2.325 \times 10^7 M_\odot$. The velocity and the angular velocity units are 10 km/s and 10 km/s/kpc respectively while G is equal to unity. Therefore, the energy unit (per unit mass) is $100 (\text{km/s})^2$. In these units, we use the values $c_n = 0.01$ kpc, $\alpha = 1.5$ kpc, $h = 0.02$ kpc, $M_d = 1000 - M_n$.

2. Results for the time independent model

In order to study properties of motion and the general dynamical behavior of the system, we shall use the two dimensional Hamiltonian,

$$H = \frac{1}{2}(p_r^2 + p_z^2) + V_{\text{eff}}(r, z) = E. \quad (2)$$

Here p_r, p_z are the momenta, per unit mass, conjugate to r and z , E is the star's energy while

$$V_{\text{eff}} = \frac{L_z^2}{2r^2} + V(r, z) \quad (3)$$

is the effective potential, L_z being the conserved component of the angular momentum.

The study is based on the numerical integration of the equations of motion

$$\ddot{r} = -\frac{\partial V_{\text{eff}}}{\partial r}, \quad \ddot{z} = -\frac{\partial V_{\text{eff}}}{\partial z}. \quad (4)$$

In order to distinguish between regular and chaotic motion in the time independent model we use the $r - p_r$ ($z = 0, p_z > 0$) Poincare phase plane. Here we must make clear that with the term the “system shows chaotic orbits” we mean that, for the given values of the parameters, there is a part of the surface of section covered by chaotic orbits. Orbits were started at the plane of symmetry $z = 0$, with radial and vertical velocities ≤ 30 km/s.

First, we shall present a relationship between L_{zc} and M_n , where L_{zc} is the critical value of the angular momentum (that is the maximum value of the angular momentum for which chaotic regions are observed in the $r - p_r$ phase plane for a given value of

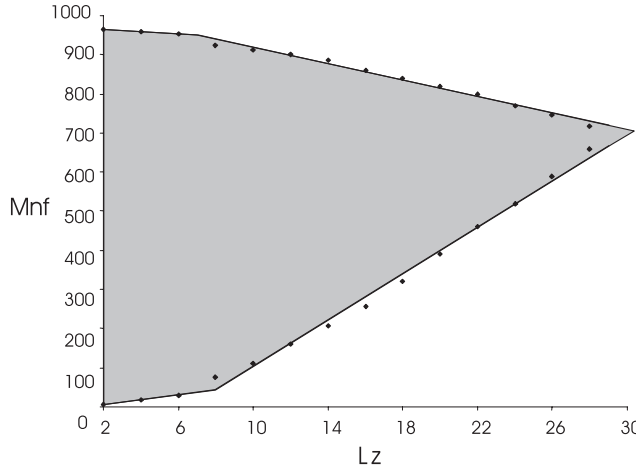


Figure 1. Relationship between L_{zc} and M_n . The values of the other parameters are given in text.

M_n , when all other parameters are kept constant). The results are shown in Fig. 1. As in the case of a disk galaxy, (see Caranicolas & Innanen 1991, Fig. 1) one observes that, a linear relationship exists between L_{zc} and M_n . What is interesting here is that, for a given value of the critical angular momentum there corresponds two values of the mass of the nucleus M_n , for which transition from regular motion to chaos occurs and not only one, as it was observed in disk galaxies. The two lines define an area in the $[L_{zc}-M_n]$ plane. For the values of the parameters within the two lines – including the lines – the $r-p_r$ phase plane shows areas of chaotic motion while, for values of the parameters outside these lines, all orbits are regular. It is interesting to observe that, as it was explained in Caranicolas & Innanen (1991), the slope of the straight lines is different for small values of L_{zc} .

Let us now explain the behavior of the system shown in Fig. 1. It is well known (see Caranicolas & Innanen 1991) that the responsibility for the appearance of chaotic orbit is the total F_z force. This force is written as:

$$F_z = -\frac{[(\alpha + (z^2 + h^2)^{1/2})^2](1000 - M_n)z}{(z^2 + h^2)^{1/2}\{[(\alpha + (z^2 + h^2)^{1/2})^2 + r^2]^{1/2}\}} - \frac{M_n z}{(r^2 + z^2 + c_n^2)^{3/2}}. \quad (5)$$

As one can see, there are two terms: the first term comes from the disk, while the responsibility for the second term is the nucleus. Figure 2 shows a plot of the two components vs. M_n for fixed values of r and z . We have taken $r = r_0 = 0.5$, $z = z_0 = 0.01$. Note that the $|F_z|$ force due to the nucleus increases as M_n increases while the other component due to the disk decreases as M_n increases. Furthermore, $|F_z|$ is proportional to M_n , while we know from Caranicolas & Innanen (1991) that $|F_z| \propto L_{zc}$. Thus, we conclude that M_n is proportional to L_{zc} , which is the result of Fig. 1.

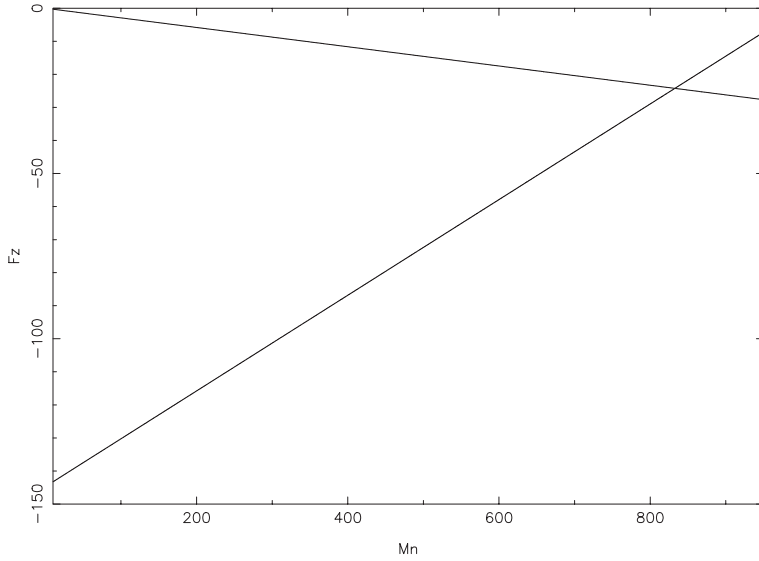


Figure 2. A plot of the two components of F_z vs. M_n for fixed values of r and z . The similarity between Figs. 1 and 2 is evident.

3. The evolving model

In this section we shall study the behavior of orbits when the system evolves with time. It is assumed that mass is transported from the disk to the nucleus. This phenomenon is described by the equations

$$M_d = M_{di} - m[1 - \exp(-kt)], \quad (6)$$

$$M_n = M_{ni} + m[1 - \exp(-kt)], \quad (7)$$

where M_{di} , M_{ni} are the initial values of the mass of the disk and nucleus, M_d , M_n are the values of mass of disk and nucleus at time t , m is the portion of the disk mass that is transported and $k > 0$ is a parameter. Note that in all cases we have $M_d + M_n = 1000$.

Again, orbits were started at the plane of symmetry $z = 0$, with radial and vertical velocities ≤ 30 km/s. In order to distinguish between regular and chaotic motion, we have calculated the maximal Lyapunov characteristic number (LCN) (see for details Lichtenberg & Lieberman 1983).

The numerical experiments show that a relationship exists between M_{nf} and L_{zc} , which is shown in Fig. 3. Here M_{nf} is the final mass of nucleus, that is, the limit of the current mass M_n as $t \rightarrow \infty$. The initial value of mass of the nucleus is $M_{ni} = 5$. One observes that the general characteristics are the same as those given in Fig. 1. Again there are two lines. For the values of the parameters in the area – including the lines – the system shows chaotic orbits, while for the values of the parameters outside the two lines all orbits are regular. Note that in this case we do not have straight lines. We believe that this is a result coming from the evolving potential, where we have an exponential evolution of the strength of the F_z force due to the exponential mass transportation.

This is justified if we look at Fig. 4, where a plot of the F_z force due to the nucleus and disk is given as a function of time t . In other words, we have replaced in (5) the value of the mass M_n with that given from equations (6) and (7). Thus, we have made the F_z force a function of time. The values of r and z are as in Fig. 2. We see that the F_z force that belongs to the nucleus increases exponentially, while that belonging to the disk decreases exponentially. Using the same arguments as in the time-independent case, we can replace F_z by M_n and t by L_z to get a similar figure to Fig. 3. Note that we do not expect to get the same results from Figs. 3 and 4. The comparison only shows that the numerically obtained results can be qualitatively reproduced using theoretical arguments together with some numerical evidence.

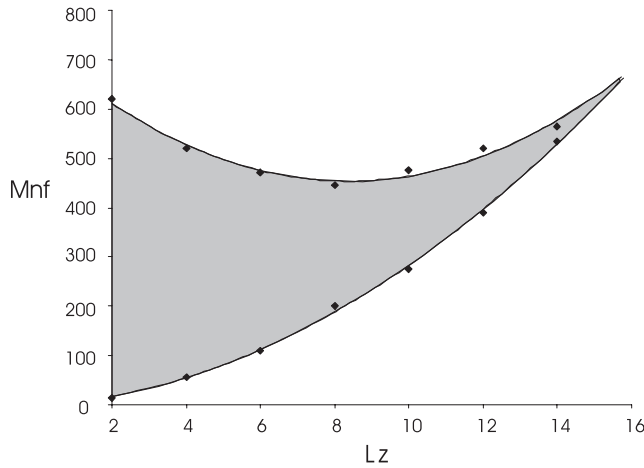


Figure 3. Same as Fig. 1 but for the evolving model. Note that the area where chaotic motion is observed is not defined by straight lines.

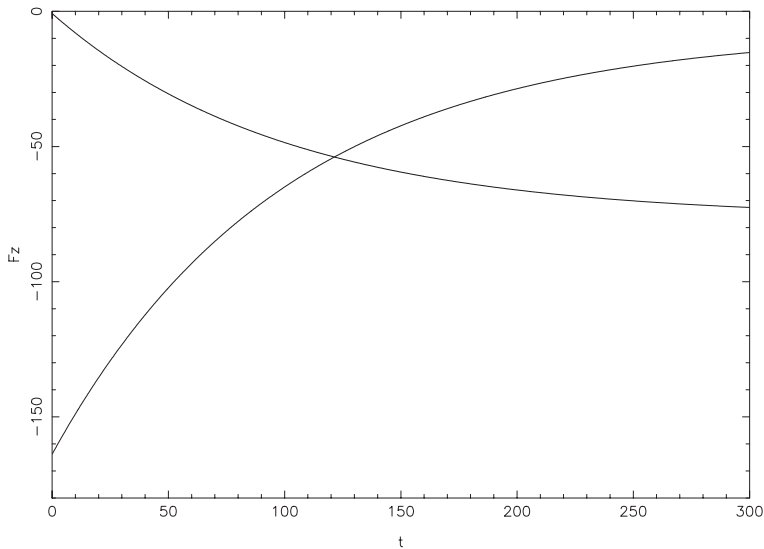


Figure 4. Same as Fig. 2 but for the evolving model. Note that the similarity to Fig. 3 is evident.

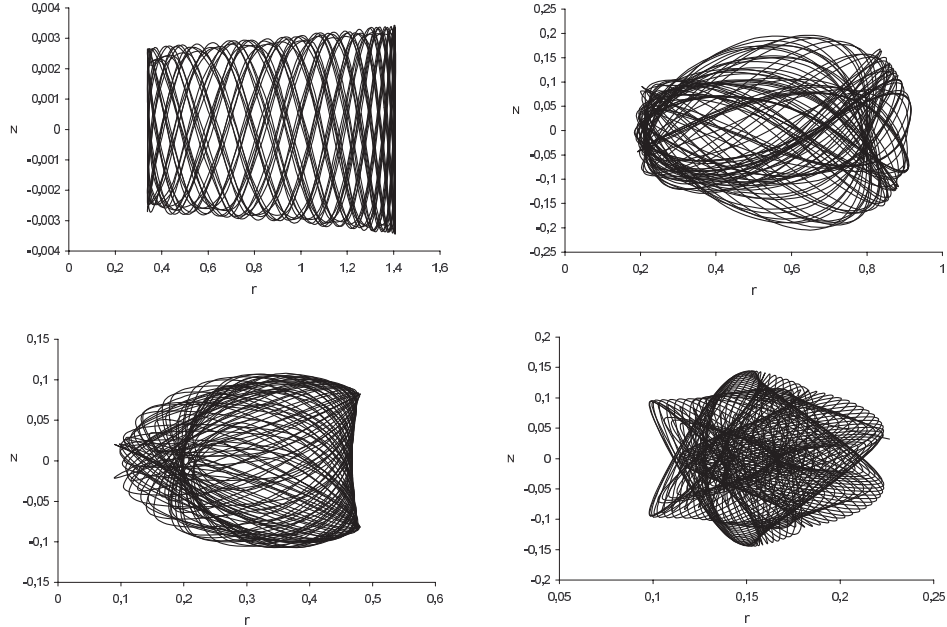


Figure 5. Evolution of an orbit in the time dependent model. It is interesting to note that the orbit switches between regularity and chaos and *vice-versa*.

It is interesting to follow the evolution of orbits in the time dependent potential. According to the diagram shown in Fig. 3, an orbit with a given value of angular momentum and a small initial value of the mass of the nucleus starts as regular, then becomes chaotic, when the mass of nucleus evolves between the limits given by the two lines in Fig. 3, and later becomes again regular. The results are shown in Fig. 5. The orbit has initial conditions $r = 1.536$, $z = 0 = p_r = 0$, while the value of $p_z \leq 30$ km/s is found using the initial value of energy (at $t = 0$). The initial value of energy, which is not conserved, is $E = -450$, $k = 0.01$ while the value of angular momentum is $L_z = 8$. The upper left shows the orbit when $5 \leq Mn \leq 55$, the upper right when $200 \leq Mn \leq 250$, down left $440 \leq Mn \leq 490$, down right $705 \leq Mn \leq 755$. Figure 6 shows the LCN of the orbit shown in Fig. 5. The numbers indicate the areas of regularity 1, 4 and the areas where the orbit is chaotic 2, 3.

From all the above, one can see that, in our time-dependent quasar dynamical model, the regular or chaotic nature of orbits strongly depends on the mass of nucleus with two transition points for a given value of critical angular momentum. On the contrary, in dynamical models for disk galaxies with active nuclei things are different where for a low value of the critical angular momentum there corresponded only one value of mass of nucleus for which transition from regular to chaotic motion occurred (see Fig. 1 in Caranicolas & Papadopoulos 2003).

This significant difference is explained if we make a plot of the F_z -force vs. M_n . Figure 7 shows such a plot. In other words, the diagram in Fig. 7 is similar to that in Fig. 4 but for the disk galaxy described in Caranicolas & Papadopoulos (2003). Note that the F_z -force due to the disk is practically constant while the attractive nuclear

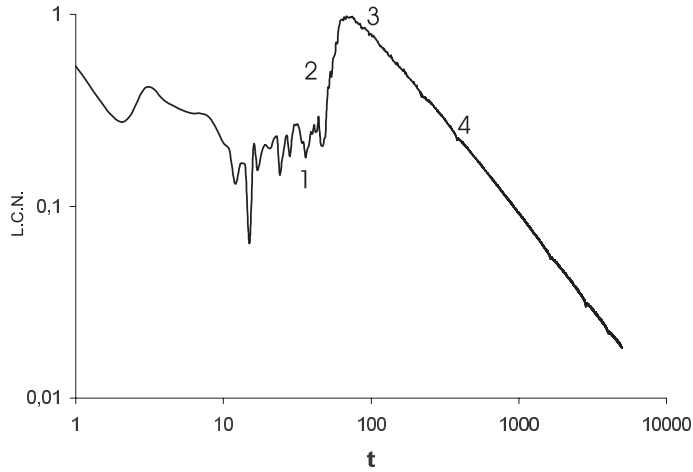


Figure 6. LCN for the orbit shown in Fig. 5. The areas of regularity are indicated as 1, 4 and the areas where the orbit is chaotic are indicated as 2, 3.

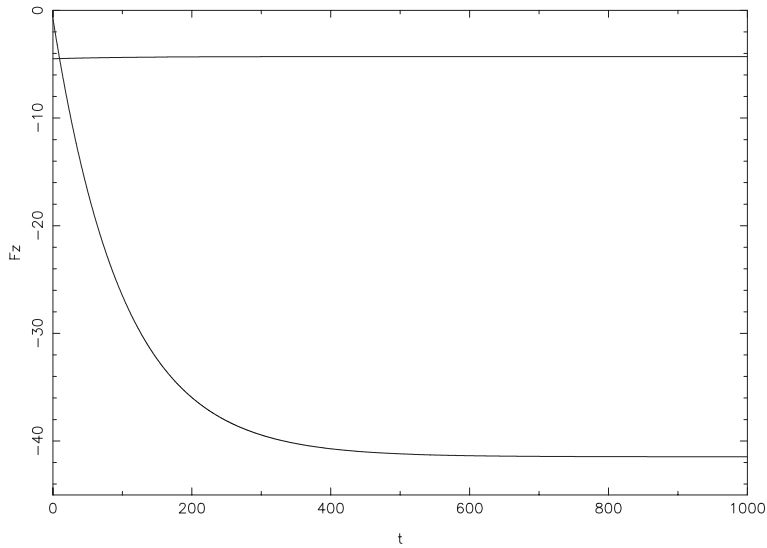


Figure 7. Same as Fig. 2 but for the disk-galaxy evolving model described in Caranicas and Papadopoulos (2003). Details are given in text.

F_z -force increases rapidly as M_n increases. As one can see, the area defined between these two curves is meant for very small values of M_n . Therefore, it seems very possible that it is difficult to detect it through numerical experiments. Furthermore, this area is of no importance because, in any case, it is negligible.

4. Discussion

In this article we have adopted a simple axially symmetric model with a disk and a super-massive nucleus in order to describe the motion in a quasar. Using this model we

tried to derive relationships connecting the physical parameters, such as the mass of the nucleus and the angular momentum, to the ability of the system to display chaotic motion. Our choice had two targets. As our knowledge for the dynamical behavior of those systems is poor, the first target was to derive results for their dynamical properties. The second target was to compare the outcomes with the already available results for galaxies.

A linear relationship was found to exist between the mass of nucleus and the value of critical angular momentum in order to observe chaotic motion. The same relationship, but not linear, seems also to exist for the evolving model. Both relationships are explained using semi-theoretical arguments together with a plot of the F_z -force which is responsible for the appearance of chaos.

Interesting results are obtained if we follow the evolution of orbits in the time-dependent model. Our numerical work shows that as mass is falling from the disk to the nucleus, an orbit can start as regular, then become chaotic and finally end up as a regular orbit.

The dynamical behavior of quasars is different from that observed in disk galaxies. The reason is that in the case of a quasar model, the transportation of mass from disk to the nucleus affects drastically the strength of the F_z -force coming from the disk, while this force remains practically constant in the case of a disk galaxy. On the other hand, the chaotic phenomena in axially symmetric quasars are seen to be smaller than those observed in non-axially symmetric quasars as the LCN (see Papadopoulos & Caranicolas 2005) in the latter case is much larger.

This paper is focused on orbital dynamics which applies to collisionless components such as stars. The effect of gaseous matter is not discussed. But, as this is a minor point, our work can be considered as a good approximation because in most galaxies the gaseous component is less than 10% of the total galactic mass.

To end up we would like to make clear that all the results of this work were based on a simple dynamical model, which was chosen for two basic reasons: the first reason was that we considered it difficult to construct, understand and handle a complicated model, while the second reason was that we wanted to start from a simple model which had similarities with the already well-known galactic disk models. On the other hand, we believe that, using the outcomes of the present work as a starting point, we will be able to make better efforts in the future.

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Chemistry of Carbon Rich Star IRAS 15194–5115

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Abstract. We have constructed two gas-phase models to study the chemistry of circumstellar envelope surrounding the carbon-rich variable star IRAS 15194–5115. The network used consists of 3893 reactions involving 397 gas-phase species. The derived fractional abundances for many molecules are in excellent agreement with values obtained from observations. The predicted column densities from the two models go well with the observed values of carbon star IRC + 10216. The dominant formation routes for three groups of species are discussed through the inner and outer envelopes.

Key words. Molecular data—molecular processes—stars: AGB—stars: individual: IRAS 15194–5115.

1. Introduction

IRAS 15194–5115 is the third brightest carbon star at $12\ \mu\text{m}$ and the brightest one known in the southern hemisphere. The object was identified as a carbon star from optical and infrared spectroscopy (Meadows *et al.* 1987), and from its position in the [K-L, $12\ \mu\text{m}$ – $12\ \mu\text{m}$] color–color diagram (Epchtein *et al.* 1987). Le Berte & Epchtein (1990) have monitored the star in the near-infrared and derived a period of 578 days. Feast *et al.* (2003) presented JHK_L observations for this star extending over about 18 years. Based on their data, the pulsation period of IRAS 15194–5115 is 575 days and has remained essentially constant over this time span. Their results suggest that there are obscuration changes due to the formation of dust clouds of limited extent in the line of sight. From the radio and far-infrared observations given by Ryde *et al.* (1999), they suggest that the object appears to be a highly evolved AGB-star, but the carbon star properties combined with inferred low $^{12}\text{C}/^{13}\text{C}$ ratio make the evolutionary status of this star uncertain. It may have been a J-star for which the $^{12}\text{C}/^{13}\text{C}$ ratio has remained low, or it may be a star of 5 to 8 solar masses, which has recently become a carbon star due to the quenching of hot bottom burning.

IRAS 15194–5115 in the infrared region shows similarities to the well-studied carbon star IRC + 10216. However, the mass-loss rate of IRAS 15194–5115 is higher than the average carbon stars. Schoier & Olofsson (2001) measured a median mass-loss rate for carbon stars of $\sim 3 \times 10^{-7} M_{\odot}/\text{yr}$. As reported by Woods *et al.* (2003,

hereafter WSNO), this suggests that IRAS 15194–5115 is going through the superwind phase evolution at the end of the AGB, and will soon eject its entire stellar mantle. Nyman *et al.* (1993, hereafter NOJBCW) have used radio observations to study the molecular species of IRAS 15194–5115 and found that circumstellar abundances, except for a few species, are comparable to those in IRC + 10216. The most striking difference between these two objects was the $^{12}\text{C}/^{13}\text{C}$ ratio, which appears to be at least five times lower in IRAS 15194–5115. A millimeter molecular line survey for IRAS 15194–5115 and six other high mass-loss carbon stars have been presented by WSNO. Their reported abundances for 24 molecular species in the seven stars agreed to within a factor of five of each other, indicating that circumstellar envelopes (CSEs) around carbon stars have similar molecular compositions. The derived abundance ratio of IRAS 15194–5115 relative to IRC + 10216 given by WSNO is higher by a factor of ~ 4 as compared with the data presented by NOJBCW. This was explained by WSNO as being due to the difference in mass loss rates of the two stars.

Circumstellar envelopes of carbon-rich stars are very rich in molecular species. More than 50 different molecular species have been observed to date in IRC + 10216. In addition to simple carbon-bearing molecules, large carbon-chain molecules like cyanopolynes, hydrocarbon radicals, carbenes, and sulphur and silicon bearing species have been detected in its envelope (Guelin *et al.* 1998; Olofsson 1997; Olofsson 1994; Cernicharo *et al.* 1991). These species can be divided into parent and daughter molecules. Parent molecules are formed in thermodynamic equilibrium in the stellar photosphere and are ejected with the outflowing material. When the ejected gas reaches the external layers of the envelope where UV photons can penetrate, parent molecules are dissociated into reactive daughter products that give rise to the outer envelope chemistry (Fuente *et al.* 1998).

Early chemical schemes of carbon rich CSEs based on models of interstellar chemistry were published by Glassgold *et al.* (1986), Nejad & Millar (1987) and Glassgold *et al.* (1987). These types of models are in reasonably good agreement with observations, but they could not explain the generation of the larger molecules. Recently, several new models have been proposed for predicting the larger molecules for the brightest carbon rich star IRC + 10216. Millar *et al.* (2000), developed a new chemical model of the outer CSE of IRC + 10216 in which carbon-bearing molecules up to 23 carbon atoms were included. This work was an extension of the earlier work of Bettens & Herbst (1995). In their new model, they extended standard models of gas-phase interstellar chemistry to produce unsaturated molecules as large as fullerenes. The chemical network used in the previous model have been used to model the enhanced density shells around IRC + 10216 (Brown & Millar 2003) and the synthesis of benzene in the protoplanetary nebula (PPN) CRL618 (Woods *et al.* 2002).

Since the chemistry of IRAS 15194–5115 did not study theoretically up till now, the goal of the present paper is to study the chemistry of this object. For this purpose, we have developed two gas-phase models. The first model (model 1) is constructed to compare with the observations reported by NOJBCW, while the second model (model 2) is to compare with the recent observations of WSNO. The purpose of these two models is to predict the fractional abundance and some physical parameters observed for this object. In section 2, we present the initial parameters of the two models, equations used to calculate gas density, temperature and visual extinction and chemical network

Table 1. Initial parameters for models 1 and 2.

	Model 1	Model 2
Mass loss rate	$5 \times 10^{-5} M_{\odot}/\text{yr}$ (NOJBCW)	$1 \times 10^{-5} M_{\odot}/\text{yr}$ (WSNO)
Expansion velocity	22 km/s (NOJBCW)	21.5 km/s (WSNO)
Initial radius	5×10^{15} cm	
Final radius	5×10^{17} cm	
UV radiation	Standard interstellar value	
Ionization rate (ζ)	Standard interstellar value	
H ₂	See equation (1)	
C/O	1.19	

database. In section 3, we show the results of the two models and discuss the chemistry of several species in inner and outer parts of the envelope. Conclusions are given in section 4.

2. Physical models

For this study, we have assumed that the gas in IRAS 15194–5115 expands in a spherically symmetric outflow with constant velocity. We have adopted the expansion velocity and mass-loss rate of (NOJBCW) for model 1 and (WSNO) for model 2 (see Table 1). The gas density of H₂ is estimated as

$$n_{H_2}(r) = \frac{\dot{M}}{4\pi r^2 v_e m_H}, \quad (1)$$

where $\dot{M}$ is the stellar mass loss rate, v_e is the expansion velocity, r is the distance from the star and m_H is the mass of molecular hydrogen (Willacy & Millar 1997). The temperature is given as a function of the shell radius according to the following temperature profile

$$T(r) = \max \left[100 \left(\frac{r}{r_i} \right)^{-0.79}, 10 \right], \quad (2)$$

where r_i is the initial radius of the shell (Millar & Herbst 1994). The visual extinction, A_v , is estimated according to the relation

$$\frac{N_{H_2}}{A_v} = 10^{21} \text{ cm}^{-2} \text{ mag}^{-1}, \quad (3)$$

where the column density of H₂ is given by

$$N_{H_2} = n_{H_2}(r)r. \quad (4)$$

Table 2. Adopted initial fractional abundances of parent species with respect to $n(\text{H}_2)$.

Species	Initial fractional abundance
He	0.15
CO	$6.0\text{E} - 4$
N_2	$2.0\text{E} - 4$
C_2H_2	$2.0\text{E} - 5$
HCN	$1.0\text{E} - 5$
CS	$1.0\text{E} - 6$
CH_4	$2.0\text{E} - 6$
H_2S	$1.0\text{E} - 6$
SiH_4	$4.0\text{E} - 6$
SiS	$3.0\text{E} - 7$

The calculations of the radial distribution of molecular abundances start at a radius of 5×10^{15} cm, until the outer radius of 5×10^{17} cm. According to equation (2), the temperature varies from ~ 100 K at the inner radius to 10 K at the outer radius in both models. Most of the predicted species give their peak values between 9×10^{15} cm and 9×10^{16} cm, where the temperature varies from ~ 60 to 10 K. This range of temperature roughly agrees with the observed kinetic temperature range of ~ 30 to 10 K given by WSNO and that given in Fig. 4 of Ryde *et al.* (1999). With respect to the density derived from equation (1) in the range 9×10^{15} cm to 9×10^{16} cm, models 1 and 2 give density range of $\sim 8.0 \times 10^4$ to $8 \times 10^2 \text{ cm}^{-2}$ and $\sim 5.0 \times 10^5$ to $5 \times 10^3 \text{ cm}^{-2}$ respectively. These values are compatible with the observed density range $\sim 10^5$ to 10^2 cm^{-2} , given by WSNO.

Initial parameters like mass loss, expansion velocity, UV radiation and cosmic ray ionization rate are given in Table 1 for our two models. Initial fractional abundances of parent species with respect to $n(\text{H}_2)$ are given in Table 2. These values are typical of carbon-rich AGB stars. We take into account the effect of self-shielding for H_2 and CO molecules against the photo-dissociation. We have used the chemical reaction network and rate coefficients (RATE 95 data base) of Millar *et al.* (1997). The chemical network has 3893 gas-phase reactions and 397 species.

3. Results and discussion

3.1 Comparison between observations and calculations

The results of model 1 show excellent agreement with the observation given by NOJBCW. Table 3 displays a comparison between calculated and observed fractional abundances. The calculated abundances of all species, except for C_3N and C_4H , are in the same order of magnitude as the observed values. The ratios between calculated

Table 3. A comparison between calculated (model 1) and observed (Nyman *et al.* 1993) fractional abundances of IRAS 15194–5115.

Species	NOJBCW	Model 1
CN	1.70E – 06	1.90E – 06
CS	1.20E – 06	1.00E – 06
SiO	4.80E – 07	3.00E – 07
SiS	9.70E – 07	9.80E – 07
HCN	1.00E – 05	9.90E – 06
HNC	9.50E – 08	5.40E – 08
SiC ₂	9.00E – 07	4.00E – 07
C ₂ H	1.60E – 05	1.60E – 05
C ₃ H	3.60E – 08	3.70E – 08
C ₄ H	7.40E – 06	1.30E – 07
C ₃ H ₂	9.60E – 08	4.80E – 08
C ₃ N	1.70E – 07	7.10E – 10
HC ₃ N	6.40E – 07	2.50E – 07

and observed abundances for most species vary from ~ 0.4 to 1, while abundances for C₃N and C₄H are two orders less than the observed ones.

The results of model 2 are compared with the recent observations given by WSNO. The results given in Table 4, agree well with the observations. The predicted fractional abundances agree within 1–2 orders of magnitude with the observed values. CN, CS, SiO, SiS, HCN, HNC, C₂H, C₃H, C₃H₂ and HC₃N indicate good agreement with similar observed species. The ratios between calculated and observed values for the previous species vary from 0.1 to 1. The observed species C₃N, HC₅N, C₂S, C₃S and CH₃CN are given as the upper limit, therefore our results of these species could possibly be reasonable. SiC₂ and C₄H the only two species show less agreement with observations. Their abundances are up to two orders of magnitude less than the observations. In general, the obtained predictions using model 1 match more closely the observations than model 2.

Table 5 shows good agreement between the column densities derived from our two models and the observations of IRC + 10216. We compared our results with that of IRC + 10216, because there are no observed column densities published for the species of IRAS 15194–5115. The first column of Table 5 is the observed column densities given by WSNO, while the second column shows the observations collected from other sources. The calculated column densities for most species given in Table 5 fit well with the observed ones, except for the following species: C₂H which is one order of magnitude larger than the observed values, C₃N is about one order less than the observed value and the large molecules HC₅N, HC₇N and HC₉N are 1–3 orders less than the observed values. The large value of C₂H in our object comparing with IRC + 10216, could be due to the observed fractional abundance of this

Table 4. A comparison between calculated (model 2) and observed (Woods *et al.* 2003) fractional abundances of IRAS 15194–5115.

Species	WSNO	Model 2
CN	1.20E – 06	2.50E – 06
CS	2.00E – 06	9.90E – 07
SiO	1.70E – 06	3.00E – 07
SiS	4.60E – 06	1.00E – 06
HCN	*1.20E – 05	9.90E – 06
HNC	3.30E – 07	4.90E – 08
SiC ₂	8.10E – 06	5.60E – 07
C ₂ H	1.50E – 05	2.00E – 05
C ₃ H	9.20E – 08	6.90E – 08
C ₄ H	2.80E – 05	8.20E – 07
C ₃ H ₂	6.00E – 07	3.40E – 07
C ₃ N	< 4.80E – 07	6.70E – 09
HC ₃ N	2.10E – 06	4.90E – 07
HC ₅ N	< 4.50E – 06	1.50E – 08
C ₂ S	< 3.00E – 08	4.20E – 09
C ₃ S	< 7.20E – 08	4.80E – 09
CH ₃ CN	< 1.20E – 06	1.30E – 09

*HCN = $5.5 \times \text{H}^{13}\text{CN}$.

molecule in our object which is about 5 times that of IRC + 10216 (see Table 9 of WSNO).

Bachiller *et al.* (1997, Fig. 5) reported that the abundances of CN, HNC and HCO⁺ increase dramatically when the star transits from the stage of AGB, through the proto-PN and PN phases. CN/HCN ratio has a wide varying range from 0.07–0.68 for high mass-loss AGB stars (WSNO), while it has very narrow range ~ 0.65 –0.7 in low mass-loss stars (Olofsson *et al.* 1993). Our derived ratio of CN/HCN from both models fit well with the observations of NOJBCW and WSNO. The observed values of CN/HCN given by WSNO and NOJBCW or our derived values from models 1 and 2 indicate that IRAS 15194–5115 is an early type carbon star (Table 6). Our predicted HNC/HCN ratio also fits well with the observations of NOJBCW and Omont (1993). This value increases by two orders of magnitude in PN phase as compared to CSE. The molecular ion HCO⁺ has not been observed in our object, but its derived column density given in Table 5 fits well with the observed values of IRC + 10216. As given in Table 6, the ratio of HCO⁺ relative to HCN matches the observed value given by Omont (1993). The previous three ratios, in addition to the ¹²CO/¹³CO ratio, can be used to estimate the evolutionary status of carbon stars. As the star evolves along AGB, the ratio ¹²CO/¹³CO, which depends on stellar mass, is expected to increase gradually from an initial low value of ~ 10 –15 (Abia *et al.* 2001).

Table 5. A comparison between observed (IRC + 10216) and calculated (IRAS 15194–5115) column densities (cm^{-2}).

Species	Observations		Predictions	
	WSNO	Other sources	Model 1	Model 2
CN	8.3E + 14		2.74E + 15	1.22E + 15
CS			7.80E + 15	1.93E + 15
HCN			7.72E + 16	1.93E + 16
HNC	2.0E + 13		7.78E + 13	2.38E + 13
SiO			2.34E + 15	5.85E + 14
SiS			7.64E + 15	1.46E + 15
SiC ₂			7.92E + 13	1.62E + 14
C ₂ H	8.9E + 14	3 – 5E + 15(1, 2, 3)	2.30E + 16	3.24E + 16
C ₃ H	2.1E + 13	3E + 13(1, 4)	5.33E + 13	1.12E + 14
C ₄ H	9.3E + 14	2 – 9E + 15(1, 3, 4, 5)	1.87E + 14	1.20E + 15
C ₅ H		2 – 50E13(1, 4)	1.58E + 13	5.54E + 13
C ₆ H		3 – 30E13(1, 4)	3.89E + 12	2.95E + 13
C ₃ H ₂	1.2E + 13	2E + 13(4)	3.53E + 13	1.65E + 14
C ₃ N	2.0E + 14	2 – 4E14(4, 6)	4.93E + 11	3.26E + 12
C ₅ N		3E + 12(6)	1.58E + 13	1.58E + 12
HC ₃ N	9.1E + 14	1 – 2E15(4, 6)	4.95E + 14	7.15E + 14
HC ₅ N	5.8E + 15	2 – 3E14(4, 6)	7.78E + 12	2.19E + 13
HC ₇ N		1E14(4)	6.12E + 10	3.82E + 11
HC ₉ N		3E + 13(4)	1.73E + 09	1.15E + 10
HCO ⁺		3E + 12(7)	1.28E + 13	1.66E + 12
CH ₃ CN	9.9E + 12	6E + 12(8)	3.60E + 12	9.36E + 11

References: (1) Guelin *et al.* (1997); (2) Groesbeck *et al.* (1994); (3) Avery *et al.* (1992); (4) Kawaguchi *et al.* (1995); (5) Dayal & Bieging (1993); (6) Guelin *et al.* (1998); (7) Olofsson (1997); (8) Guelin & Cernicharo (1991).

Table 6. A comparison between observed and calculated ratios of CN, HNC and HCO⁺ relative to HCN.

	NOJBCW	WSNO	Omont 1993	Model 1	Model 2
CN/HCN	0.17	0.1	0.075	0.19	0.25
HNC/HCN	9.5E – 03	2.7E – 02	4.0E – 3	5.5E – 03	4.9E – 03
HCO ⁺ /HCN	–	–	< 2.0E – 4	8.9E – 4	2.3E – 4

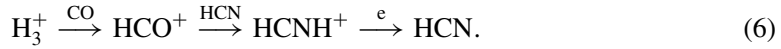
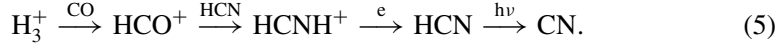
3.2 Molecular formation

The chemistry through the circumstellar envelope was thought to be consisting of three stages: the inner chemistry, where the parent molecules begin to dissociate; the middle chemistry, where the daughter molecules reach their peak abundances and the

outer chemistry, where interstellar photons dissociate all molecules (Willacy & Millar 1997). Shock waves were thought to affect the dynamics and chemistry of the inner layers of CSE (Kerschbaum & Olofsson 2001). A dynamical and chemical model was established for the inner wind of carbon-rich star IRC + 10216 by Willacy & Cherchneff (1998). This model takes the effect of pulsation-driven shocks on the gas envelope close to the stellar photosphere.

3.2.1 *CN, HCN and HNC species*

CN is the daughter product of the parent molecule HCN. Its formation occurs due to the photo-dissociation of HCN only in the inner envelope and both HCN and HNC in the outer envelopes. The parent molecular ion HCNH^+ is responsible for the formation of HCN and HNC via dissociative recombination with electrons in the inner and outer envelopes. The main formation routes of CN, HCN and HNC in the inner part are initiated by reactions involving H_3^+ .



In the outer envelope the dominant formation routes for the previous three species are initiated by carbon ion C^+



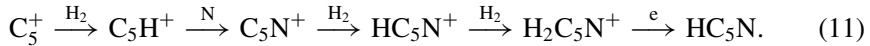
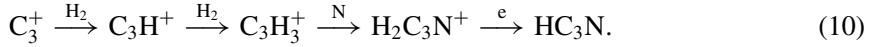
CN radial abundance distribution shows an increase and then a gradual decrease through a peak at 2×10^{16} cm (Fig. 1). HNC has the same behaviour as CN, except that its peak is at 3×10^{16} cm.

3.2.2 *Organo-nitrogen species*

The main route that produces cyanopolynes in the inner envelope occur via reactions between CN and C_{2n}H_2 , where $n = 1, 2, 3$ and 4,



In the outer region, the following main routes are responsible for the formation of cyanopolynes series:



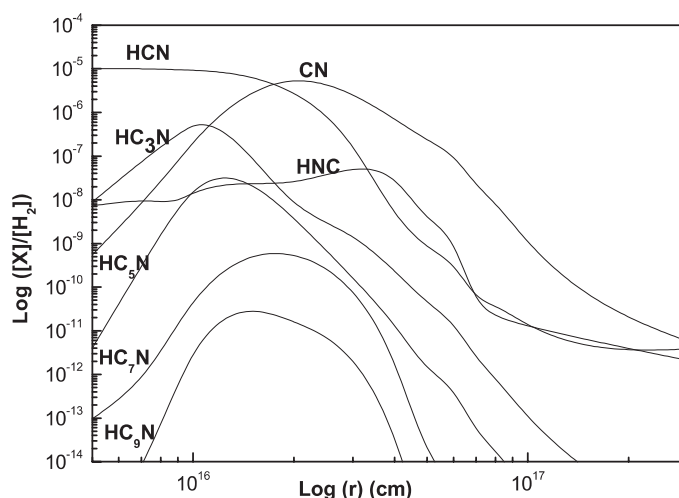
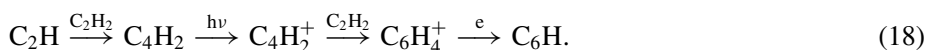
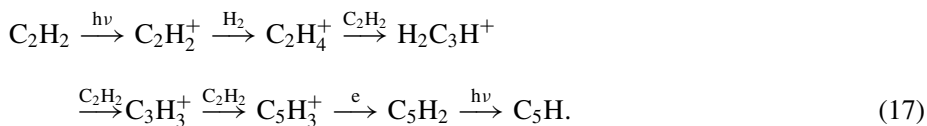
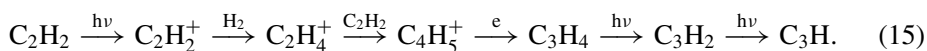


Figure 1. The radial distribution of CN, HNC and organo-nitrogen species (model 2).

Figure 1 shows the radial distribution of the fractional abundances for organo-nitrogen species HCN, HC₃N, HC₅N, HC₇N and HC₉N. HC₃N and HC₅N molecules peak at about 1×10^{16} cm, while HC₇N and HC₉N molecules roughly peak at 2×10^{16} cm.

3.2.3 Hydrocarbon species

The formation of hydrocarbon species C₂H, C₃H, C₄H, C₅H and C₆H in the inner envelope occur according to the following routes:



The route schemes 15 and 17 initiated by photo-dissociation of acetylene C₂H₂, followed by ion-neutral reactions, dissociative recombination with electrons and end by photo-dissociation. While route scheme 18 is initiated by neutral-neutral reactions of C₂H₂ with its primary daughter C₂H, followed by ion-neutral reactions and ends

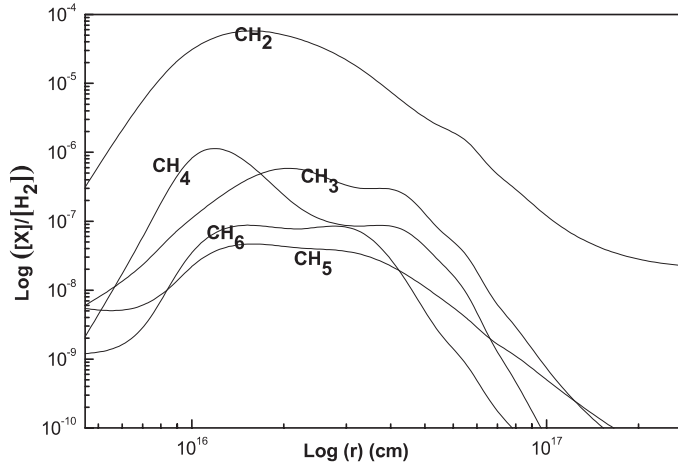
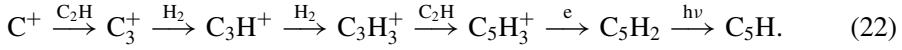
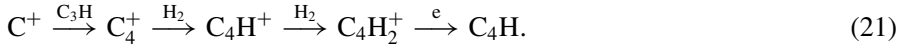
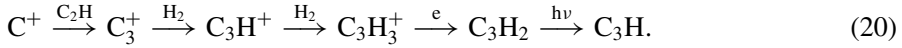


Figure 2. The radial distribution of hydrocarbon species (model 2).

by dissociative recombination with electrons. In the outer envelope, hydrocarbons are formed mainly as follows:



These routes are started by photo-dissociation of C_2H_2 or ion-neutral reactions of C^+ , followed by ion-neutral reactions of H_2 , dissociative recombination with electrons and photo-dissociation. Figure 2 displays the predicted radial fractional abundances distribution of hydrocarbons. The peak abundance of these species ranges from $\sim 9 \times 10^{15}$ – 2×10^{16} cm.

4. Conclusion

We have developed two gas-phase chemical models for the carbon star IRAS 15194–5115. These two models represent the first theoretical trial to model the chemistry of this object. The results of the theoretical models predict clearly the observed species in IRAS 15194–5115. The fractional abundances for several species are derived and compared with observed values. The column densities were also derived and compared with those from the observation of carbon star IRC + 10216. The results of both fractional abundances and column densities show very good agreement with the observations, except for very few species. In general, model 1 predicts the observation better than model 2. The ratios of CN/HCN , HNC/HCN and HCO^+/HCN are

in good agreement with observations of carbon rich stars. Finally, we presented the dominant formation routes responsible for producing CN, HNC, organo-nitrogen and hydrocarbon species in the inner and outer layers of the envelope.

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Orbital Evolution and Orbital Phase Resolved Spectroscopy of the HMXB Pulsar 4U 1538–52 with RXTE-PCA and BeppoSAX

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Abstract. We report here results from detailed timing and spectral studies of the high mass X-ray binary pulsar 4U 1538–52 over several binary periods using observations made with the Rossi X-ray Timing Explorer (RXTE) and BeppoSAX satellites. Pulse timing analysis with the 2003 RXTE data over two binary orbits confirms an eccentric orbit of the system. Combining the orbital parameters determined from this observation with the earlier measurements we did not find any evidence of orbital decay in this X-ray binary. We have carried out orbital phase resolved spectroscopy to measure changes in the spectral parameters with orbital phase, particularly the absorption column density and the iron line flux. The RXTE-PCA spectra in the 3–20 keV energy range were fitted with a power law and a high energy cut-off alongwith a Gaussian line at ~ 6.4 keV, whereas the BeppoSAX spectra needed only a power law and Gaussian emission line at ~ 6.4 keV in the restricted energy range of 0.3–10.0 keV. An absorption along the line of sight was included for both the RXTE and BeppoSAX data. The variation of the free spectral parameters over the binary orbit was investigated and we found that the variation of the column density of absorbing material in the line of sight with orbital phase is in reasonable agreement with a simple model of a spherically symmetric stellar wind from the companion star.

Key words. Stars: binaries: eclipsing—stars: pulsars: individual: 4U 1538–52—X-rays: stars.

1. Introduction

The orbits of binary systems evolve with time due to orbital angular momentum changes. This change of orbital angular momentum can happen due to mass loss from the system, mass transfer within the system and tidal interactions between the components. X-ray binary pulsars are suitable candidates to study such orbital evolution because of the following reasons. Firstly, in all these systems the normal companion

star is transferring matter to the compact star either through wind capture or through Roche-lobe overflow. Secondly, the strong gravitational pull of the compact object raises tides on the companion star which tend to circularize and synchronize the binary orbit. This causes orbital angular momentum changes which result in the change of orbital period. Thirdly, pulse timing analysis allows us to obtain an orbital ephemeris history with an accuracy upto a few seconds. Thus monitoring X-ray binary pulsars over a time span of few decades allows us to measure the rate of change of the orbital period. For most X-ray binaries the orbital ephemeris clearly deviates from the linear trend. Including a quadratic term improves the fit in several cases (Cen X-3, Nagase *et al.* 1992; LMC X-4, Levine *et al.* 2000, Naik & Paul 2004; SMC X-1, Wojdowski *et al.* 1998) indicating a non-vanishing orbital period derivative. The many X-ray satellite missions have thus made it possible to measure the rate of change of the orbital period, showing that for High Mass X-ray Binaries (HMXB) the orbital evolution time scales are of the order of a million years.

The X-rays emitted from the compact object in HMXBs and subsequently reprocessed in the stellar wind can be used to probe the stellar wind of the massive companion, a supergiant or a Be star. An investigation of the X-ray spectrum from the compact object at different orbital phases can give useful insight into the morphology of the stellar wind. The absorbing column density in the line of sight would vary with orbital phase and is a function of the mass-loss rate, stellar wind velocity and acceleration method and also depends on any anisotropy in the stellar wind structure. Moreover, a binary system with an elliptical orbit would be a better candidate in this type of study since it would have a larger variation of the parameters that determine the column density at different orbital phases. If the compact object is a pulsar, the orbit can often be determined precisely, which is an added advantage. Orbital phase resolved spectroscopy was carried out recently on some HMXB pulsars like X 1908 + 075 (Levine *et al.* 2004) and GX 301–2 (Mukherjee & Paul 2004) and very different results were obtained.

4U 1538–52 is an X-ray pulsar, first detected with the UHURU satellite (Giacconi *et al.* 1974). Regular X-ray pulsations with a period of 529 s were later discovered with Ariel 5 and OSO-8 observations by Davison *et al.* (1977). In addition, the OSO-8 observations revealed a clear orbital modulation of the pulsation period and also showed good evidence for an eclipse lasting 0.6 day recurring with a period of 3.75 days. The optical counterpart was found to be an early B type supergiant star with H-alpha emission lines (Parkes *et al.* 1978). Features in the optical spectrum suggested a distance to the source of ~ 5.5 kpc and the mass-loss rate for the companion was estimated to be $\sim 10^{-6} M_{\odot} \text{ yr}^{-1}$.

Makishima *et al.* (1987) using TENMA and Robba *et al.* (1992) using EXOSAT observed pulse periods of ~ 530 s with an average spin-down rate of $\dot{P} \sim 3.9 \times 10^{-9} \text{ s s}^{-1}$. Subsequently, the Burst and Transient Source Experiment (BATSE) onboard the Compton Gamma Ray Observatory (CGRO) detected a reversal of the long term spin down to spin up, probably in 1988 (Rubin *et al.* 1997), which has also been confirmed by BeppoSAX (Robba *et al.* 2001). The spectrum has an iron K fluorescence line at 6.4 keV, a cyclotron absorption line near 20 keV and is well fitted with a power law and high energy cut-off at ~ 16 keV (Clark *et al.* 1990; Robba *et al.* 2001). X-ray eclipse phenomena were investigated by Clark *et al.* (1994) with the GINGA data.

4U 1538–52, a wind fed HMXB pulsar with a compact eccentric orbit ($e \sim 0.18$) is very suitable for the study of the orbital evolution and wind structure of the companion star. In this paper, we report pulse timing measurements with a new observation with the RXTE and also the out-of-eclipse spectral evolution with orbital phase. Better spectral coverage with BeppoSAX (going down to 0.3 keV) enables us to constrain the column density values well. We consider the pulsar to be immersed in a spherically symmetric stellar wind from the companion star and compare model column density profiles with the observed profile.

2. Observations

We observed the source with the Proportional Counter Array (PCA) of RXTE from 2003-07-31 to 2003-08-07 covering the out of eclipse phases for two binary orbits. There were twenty-five observations, each corresponding to one orbit of the satellite. Each observation was of 1.5–6.0 ks duration. For spectral analysis, we also used an archival observation from RXTE spanning from 1997-01-01 to 1997-01-05, having fourteen satellite orbit data. Each observation had 1.6–10.0 ks of on-source time. On an average, three PCUs (see next paragraph) were ON for the observations in 2003 while all the five PCUs were ON for the observations made in 1997. Both these datasets from RXTE provide a comprehensive orbital coverage. We have also used archival BeppoSAX data obtained between 1998-07-29 to 1998-08-01, covering one binary orbit. There were forty segments in this observation; one for each satellite orbit, with an exposure of 1.7–4.0 ks. The motivation for using the BeppoSAX observations is to have a better spectral coverage since the lower values of column densities are measured more accurately with BeppoSAX, which has a good response at low energies.

RXTE consists of a large-area proportional counter array (PCA), with five Xenon proportional counter units (PCUs) and sensitive in the energy range of 2–60 keV with an effective area of 6250 cm² (Jahoda *et al.* 1996). It also consists of a high energy crystal scintillation experiment (HEXTE; 15–200 keV; 1600 cm² area), and a continuously scanning all-sky monitor (ASM; 2–10 keV; 90 cm²). The RXTE–ASM long term light curve of 4U 1538–52 folded at the orbital period is shown in Fig. 1 which clearly shows the eclipse with a slow ingress and a sharp egress. The BeppoSAX observatory has a set of Narrow Field Instruments (NFI) comprising one Low Energy Concentrator Spectrometer (LECS; 0.1–10 keV, 22 cm² @ 0.28 keV), three Medium Energy Concentrator Spectrometers (MECS; 1.3–10 keV, 150 cm² @ 6 keV), one High Pressure Gas Scintillator Proportional Counter (HPGSPC; 4–120 keV, 240 cm² @ 30 keV), and a Phoswich Detection System (PDS; 15–300 keV, 600 cm² @ 80 keV). A detailed description of BeppoSAX can be found in Boella *et al.* (1997).

3. Timing analysis and results

Only the 2003 RXTE-PCA data were used for a timing analysis. The light curve was extracted from event mode data with time resolution of 0.125 s. Only photons detected in the top two layers were used. The photon arrival times were then corrected to the solar system barycenter. Figure 2 shows the complete barycentered light curve for the August 2003 observation with RXTE-PCA.

The light curve shows variability at time scales other than the regular 526.85 s pulses from the neutron star. The pulsed emission is strong enough for individual pulses to

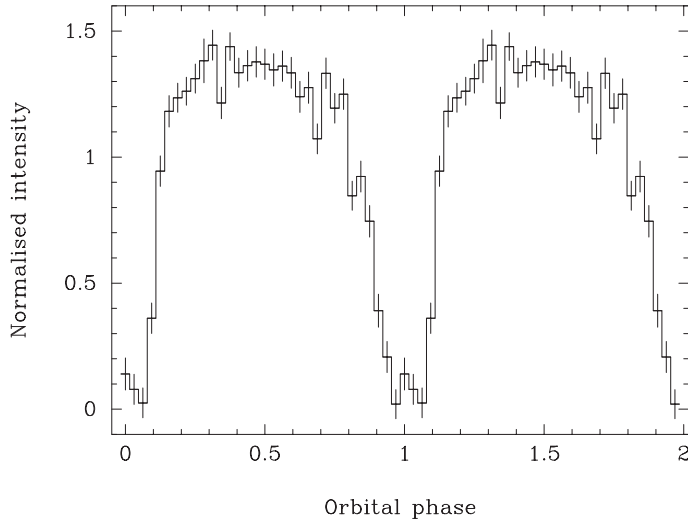


Figure 1. The RXTE-ASM light curve folded at the orbital period of ~ 3.73 days. Two orbital cycles are shown for clarity.

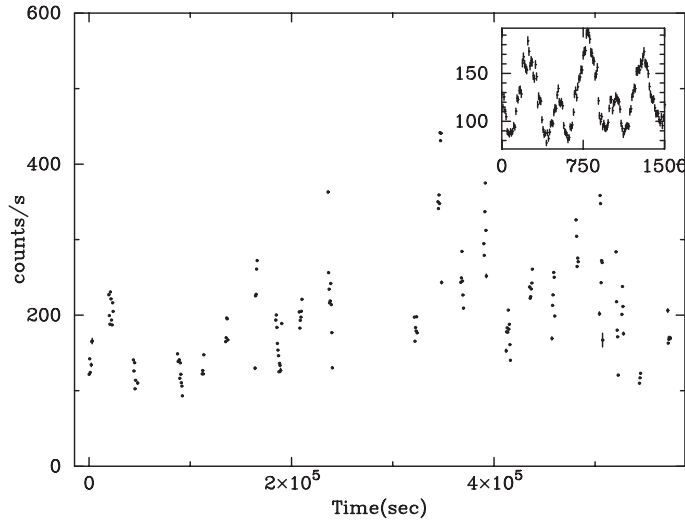


Figure 2. This plot shows the complete barycentered light curve for August 2003 data with a time bin of 526.85 s, the same as the pulse period. The inset is an expanded view of a part of the total light curve with a time bin size of 10 s. Individual double peaked pulses can be clearly seen.

be seen in the energy range from 1 keV to 40 keV. The pulse profile is double peaked with a main pulse and an inter pulse which is about half the amplitude of the main pulse. The average pulse period of the source was obtained by pulse folding and χ^2 maximisation method.

The light curve was then folded with the average spin period derived to get pulse profiles at different orbital phases. Each pulse profile is an average of two successive

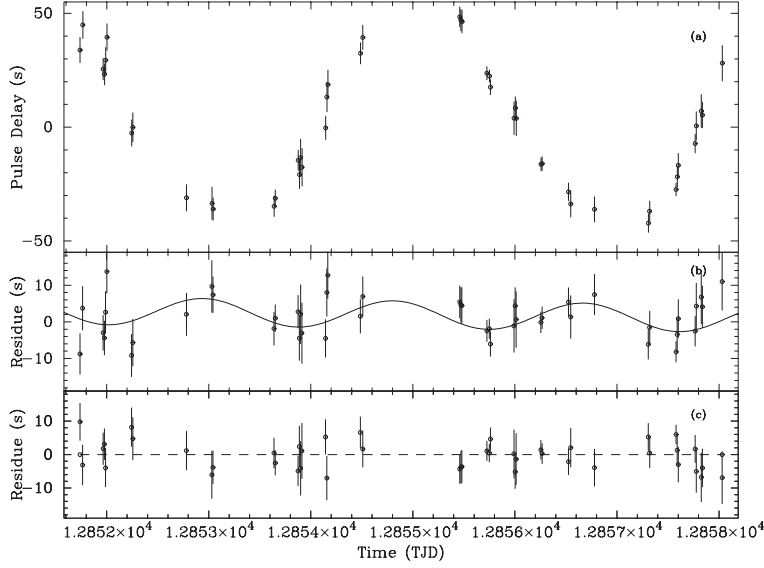


Figure 3. (a) Pulse arrival time delays measured from the RXTE-PCA observation in 2003, (b) residue of a circular fit to pulse arrival times, and (c) residue of an elliptical fit to pulse arrival times are shown here.

pulses binned into 256 pulse phase bins. A total of fortysix such pulse profiles were generated, spanned over two binary orbits of 4U 1538–52. To get the arrival times of these pulses, we decomposed each pulse profile into 128 Fourier components. The phase of the first Fourier component which also has the highest amplitude was used to find the arrival time delay of the respective pulses. Figure 3(a) shows the arrival time delay curve.

The arrival time delay curve when fitted with a circular orbital delay curve (i.e., $e = 0$) leaves out systematic residues as shown in Fig. 3(b) indicating an elliptical orbit. Subsequently, we fitted the pulse arrival time delay with a function appropriate for elliptical orbit given below:

$$\Delta t = \frac{a_x \sin i}{c} (1 - e^2) \frac{\sin(\nu + \omega)}{1 + e \cos \nu}, \quad (1)$$

$$\tan \frac{\nu}{2} = \sqrt{\frac{1+e}{1-e}} \tan \frac{E}{2}, \quad (2)$$

$$E - e \sin E = \frac{2\pi}{P_{\text{orb}}} \left(t - T_0 - \frac{P_{\text{orb}}}{2\pi} \omega \right), \quad (3)$$

$$T_0 = T_{\pi/2} - \frac{P_{\text{orb}}}{4}. \quad (4)$$

It was seen that the errors in the individual measurements of pulse arrival times were greater than the error due to photon counting statistics, possibly due to some aperiodic intensity variation over a range of time scales. While fitting the arrival delay curve to an elliptical binary orbit, the statistical uncertainties on the arrival time measurements

Table 1. The spin and orbital parameters of 4U 1538–52 with 1σ errors.

Parameter	Value
P_{spin}	526.849 ± 0.003 s
$a_x \sin i$	53.1 ± 1.5 lt-s
e	0.18 ± 0.01
ω	$40^\circ \pm 12$
$T_{\pi/2}$	52851.33 ± 0.01 MJD
a_o	47221.463 ± 0.012 lt-s
P_{orb}	3.728382 ± 0.000011 d
$T_{\pi/2}(N) = a_0 + P_{\text{orb}}N.$	

were multiplied by a constant factor to obtain a reduced χ^2 value of 1.0. This method was employed so that one can compare the orbital parameter estimations with the earlier measurements (Clark 2000). For a circular orbit, the rescaled uncertainties gives a best fit with a χ^2 value of 64 for 43 degrees of freedom. Comparing this with the reduced χ^2 value of 1 for an elliptical orbit as described above, we reconfirm an elliptical orbit for this system. The value of the five free parameters for an elliptical fit, namely P_{spin} , $a_x \sin i$, e , ω , and $T_{\pi/2}$ are listed in Table 1. We combined the new epoch measured from the RXTE data taken in 2003 with the earlier measurements listed in Clark (2000). The mid-eclipse times derived from light curves obtained using the UHURU and ARIAL V were not included because the X-ray eclipse is asymmetric as seen in the folded ASM light curve of Fig. 1. As the X-ray source goes into eclipse gradually and comes out of the eclipse rather sharply, the mid-eclipse time determined from the light curve only is likely to be inaccurate and have some systematic errors. Fitting the remaining 9 mid-eclipse times with a linear function gives a reduced χ^2 of 1.3 for 7 degrees of freedom. The residuals from a linear fit to the mid-eclipse times are shown in Fig. 4. A quadratic fit improves the reduced χ^2 by a mere 0.01 (for 6 degrees of freedom). If the mid-eclipse time record is fitted with an orbital evolution corresponding to the value reported in Clark (2000), a reduced χ^2 of 2.5 is obtained, which is larger than the same for a linear fit. We, therefore, conclude that there is no evidence for an orbital evolution in 4U 1538–52 from the available data.

4. Spectral analysis and results

For the RXTE data, we took the Standard 2 data products of PCA and extracted the source spectra for each data set using the tool *saextrect* v 4.2d for the appropriate good time intervals. Background data files were generated with the tool *runpcabackest* using background models appropriate for the source brightness and the epoch of the RXTE observation, as provided by the PCA calibration team. The background subtracted source spectra were analyzed with the spectral analysis package XSPEC v 11.2.0 (Shafer *et al.* 1989). Model spectra were convolved with the detector response matrix and fitted to the observed pulse phase averaged spectrum in the spectral range of 3–20 keV and the important spectral parameters were determined. The spectral model

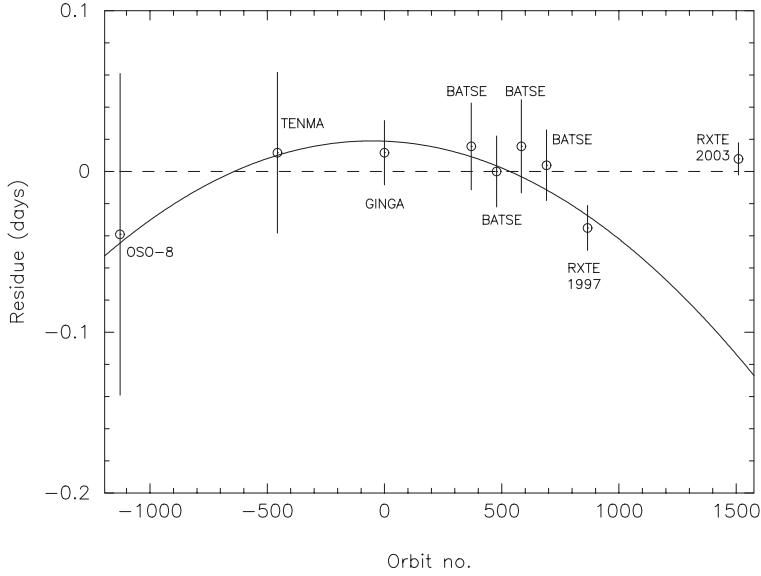


Figure 4. Differences between measured mid-eclipse times and a linear fit to the same are shown here. Solid line indicates the quadratic fit of mid-eclipse times till the 1997 observation (Clark 2000).

used was a powerlaw alongwith a high energy cut off and a Gaussian line with centre energy ~ 6.4 keV with a photoelectric absorption using Morrison and McCammon cross-sections (Morrison & McCammon 1983). All the parameters were kept free in the 2003 data set while for the 1997 data set, the line centre energy and the width were kept frozen at 6.4 keV and 0.01 keV respectively.

The source and background spectra of BeppoSAX were extracted from the MECS and LECS detectors using circular regions of radius $4'$ and $8'$ respectively. The energy range chosen for MECS was (1.8–10.0) keV while that for LECS was (0.3–4.5) keV, in which the respective instruments have large effective area and the spectral responses are well understood. For all the three observations, the MECS and LECS phase averaged out-of-eclipse spectra were fitted simultaneously with the relative normalization of the two instruments allowed to vary. All the spectra were suitably binned to allow use of χ^2 statistics. Both the MECS and LECS data of BeppoSAX were simultaneously fitted using XSPEC v 11.2.0. Due to the restricted energy range (1.8–10.0 keV), no high energy cutoff was required. The out of eclipse pulse phase averaged spectrum was first fitted with a simple model comprising of a powerlaw and a Gaussian line alongwith the photoelectric absorption using Morrison and McCammon cross-sections. The values of the line centre energy and line width as obtained in this fit were then kept frozen to fit spectra at different orbital phases.

The RXTE-PCA and the BeppoSAX spectra were all well fitted with a relatively simple model. Robba *et al.* (2001) had detected marginal evidence of a soft excess at ~ 1 keV which was resolved by modeling the spectrum with one additional component, either as a black-body of temperature 0.08 keV or a bremsstrahlung of temperature 0.1 keV. We also found that the inclusion of a soft black-body component in the phase averaged out-of-eclipse spectrum from BeppoSAX marginally improves

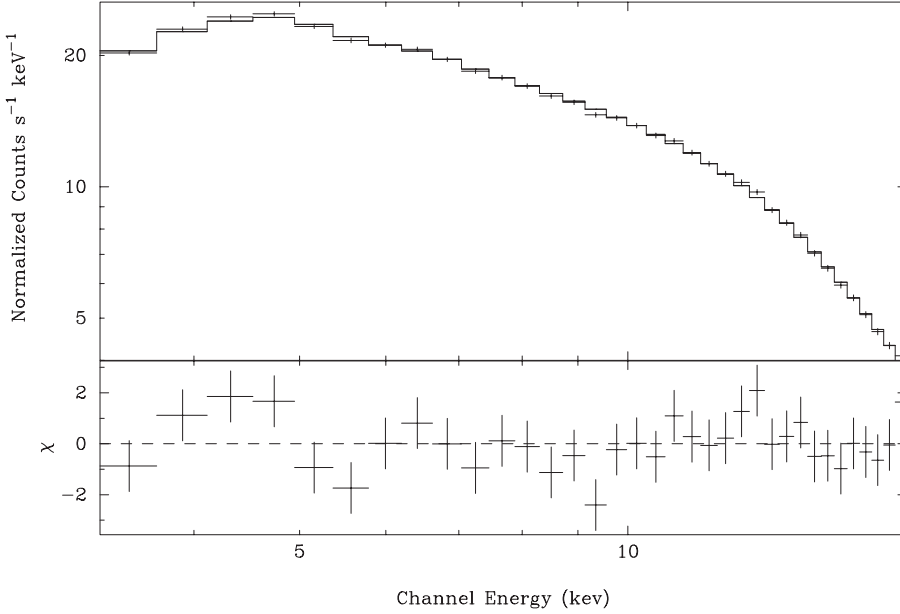


Figure 5. RXTE-PCA spectrum from an observation in 2003 corresponding to the orbital phase ~ 0.28 . The residuals are shown in the lower panel. Soft excess was not included as it lies outside the energy range.

the fit. However, for orbital phase resolved spectroscopy, we have not included a soft component as it lies outside the range of the RXTE spectrum and in the orbital phase resolved spectrum with BeppoSAX, it will be too faint to constrain. In total, there were seventy nine spectra to be fitted and with the above-mentioned model, they gave reasonably reduced χ^2 values in the range of 0.8 to 1.5 for RXTE (with 27 dof), and 0.6 to 1.3 for BeppoSAX (with 63 dof). Representative spectra, one each from the RXTE-PCA and BeppoSAX are shown in Figs. 5 and 6 respectively. The RXTE-PCA spectra analysed here also does not include the energy range of the cyclotron absorption feature.

The spectral parameters obtained from the fitting of the three data sets used in our analysis were investigated for their variation against the orbital phase of the pulsar. The photon index of the power law is seen to remain more or less constant over the binary phase having values between 1.0 to 1.5. The fluorescent iron-line flux measured with the average spectrum taken over 2000–3000 s, also does not show any considerable variation along the orbit. The cut-off energy measured with RXTE-PCA has a value ~ 14 keV and the e-folding energy is ~ 7 keV. These are slightly different from the values of the cut-off energy (~ 16 keV) and e-folding energy (~ 10 keV) measured with BeppoSAX in a broad energy band (Robba *et al.* 2001). We detected a notable variation in the equivalent hydrogen column density with orbital phase as shown in Fig. 7. It shows a smooth variation over orbital phase, increasing gradually by an order of magnitude as the pulsar approaches eclipse (mid-eclipse is defined by phase zero). At orbital phases far from the eclipse, the column density has a value of $\sim 1.5 \times 10^{22}$ H atoms cm^{-2} . In Figs. 1, 5, 6, and 7, the orbital phase is measured with respect to the mid-eclipse time.

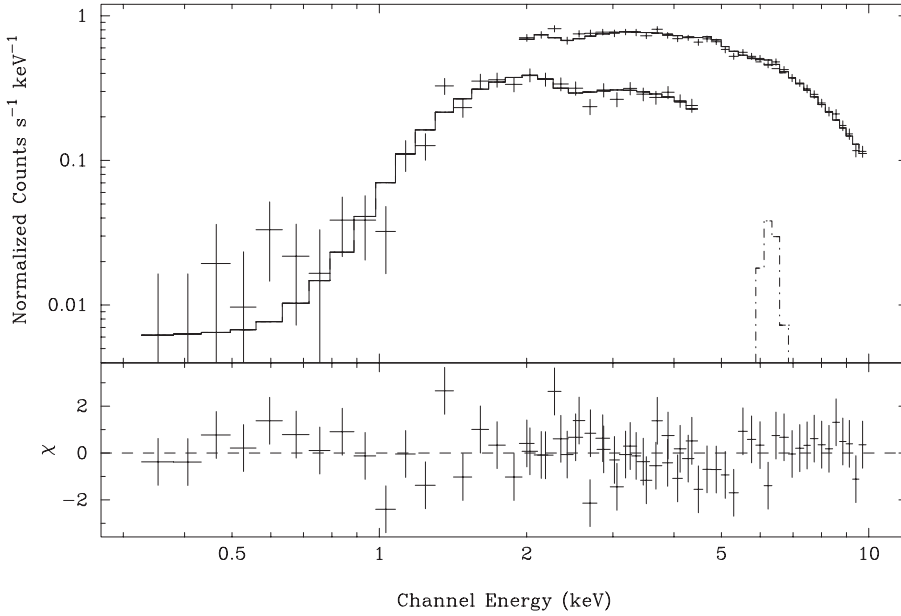


Figure 6. A BeppoSAX spectrum corresponding to orbital phase ~ 0.59 with the residuals in the lower panel. The narrow Gaussian component at 6.4 keV shown here is the iron fluorescence line.

5. Discussion

A pulse period of 526.85 s of the pulsar 4U 1538–52 determined from RXTE-PCA observation in 2003 shows that the spin-up trend of 4U 1538–52 first detected with BATSE during the 1990s has continued since then. The overall spin-up time scale ($\dot{P}/P$) during the five years of BATSE observation reported by Rubin *et al.* (1997) was $9.6 \times 10^{-12} \text{ s}^{-1}$. The average spin-up rate between the end of BATSE observations and the later observations with RXTE (1997, Clark 2000), Beppo-SAX (1998, Robba *et al.* 2001) and again with RXTE (2003, present work) is 1.4, 1.2 and $7.4 \times 10^{-11} \text{ s}^{-1}$, significantly larger than the average spin-up rate during the BATSE era. An increase in spin-up rate is expected if the overall X-ray luminosity had increased after the BATSE observations. However, it is very impractical to compare the pulsed hard X-ray flux measured by BATSE and the total X-ray flux measurements with the later instruments in a different energy band.

The value of $a_x \sin i$ that we obtained from pulse timing of the RXTE-PCA data in 2003 is $53.1 \pm 1.5 \text{ lt-s}$, which is consistent with the BATSE result and slightly smaller than that obtained from the 1997 RXTE data. We note here that the accuracy of orbital parameter determination in slow pulsars like 4U 1538–52 is considerably poorer than in fast pulsars like SMC X-1, Her X-1, etc. In the analysis of BATSE data, Rubin *et al.* (1997) considered a circular orbit whereas analysing pulse arrival times from two complete binary orbits we found the orbit of 4U 1538–52 to be eccentric with $e = 0.18 \pm 0.01$. Our result is consistent with the results obtained from a 1997 observation of this source with the RXTE-PCA (Clark 2000). The value of ω determined from the 2003 observation is smaller compared to that found by Clark from the 1997 observation.

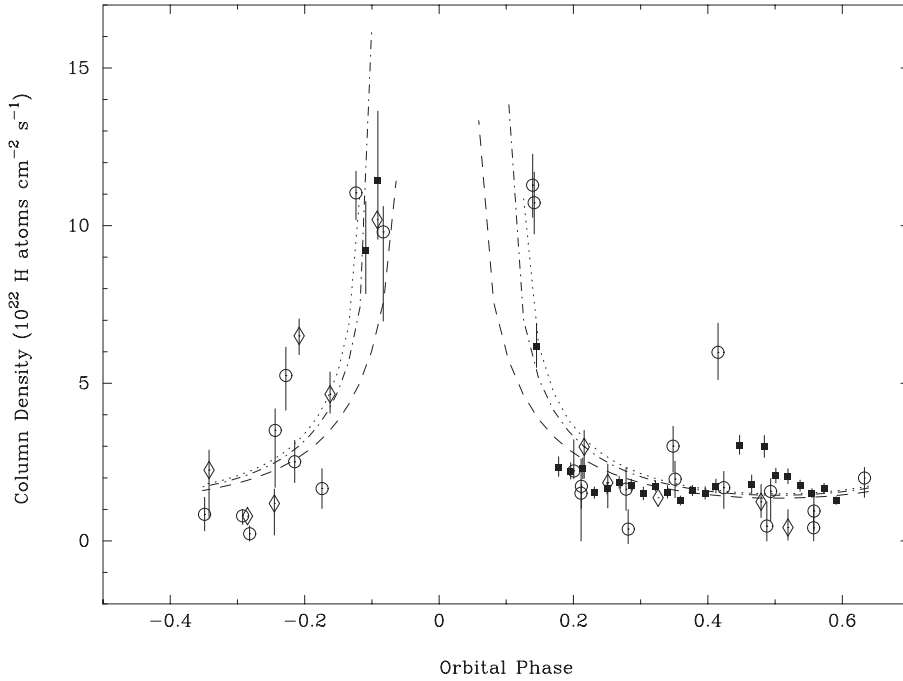


Figure 7. Variation of column density *versus* orbital phase with respect to the mid-eclipse time is shown here. The dashed line represents the model for inclination angle of 65° , the dashed-dotted line for inclination angle of 75° and the dotted line for inclination angle 85° . Diamonds, filled squares, and circles denote measurements from observations made with RXTE in 1997, BepoSAX in 1998, and RXTE in 2003 respectively. The error bars shown correspond to a 90% confidence interval.

If the difference is due to apsidal motion then the rate of apsidal motion is $\dot{\omega} = -3.8^\circ \pm 2^\circ.2 \text{ yr}^{-1}$. We found no significant evidence for orbital evolution in the system as reported from earlier measurements (Clark 2000). The value of $\dot{P}_{\text{orb}}/P_{\text{orb}}$ determined from all the available mid-eclipse times is $(2.5 \pm 1.9) \times 10^{-6} \text{ yr}^{-1}$. Combining the pulse timing results on 4U 1538–52 obtained from our observation with the earlier reported measurements, we are able to rule out an orbital decay in this system for which Clark (2000) found marginal evidence. The new upper limit presented here on possible changes of the orbital period of 4U 1538–52 supersedes previous limits considerably.¹

Tidal interaction between the two stars and mass loss from the companion star are the two main effects that cause orbital decay in the HMXB systems. The orbital evolution of several HMXB pulsar systems (Cen X-3, SMC X-1 and LMC X-4) is very well documented. The companion star mass, and the size of the binary orbit in 4U 1538–52 is very similar to the above-mentioned binaries, and therefore, the tidal effect that is expected in 4U 1538–52 should be as strong as in the three other HMXBs. In the present work, we find no clear evidence of an orbital decay in 4U 1538–52. If at all, the system shows a positive orbital period derivative, like the enigmatic system Cyg

¹After this paper was submitted, we came to know about a paper by Baykal, Inam and Beklen (2006). These authors have used a different analysis method and reached a similar conclusion about the orbital period evolution of 4U 1538–52.

X-3 (Singh *et al.* 2002). This indicates that mass loss may play a prominent role in the evolution of this system. Tidal interaction between the neutron star and the companion in close HMXBs should also circularise the orbit. The orbits of most of the close (orbital period less than 4 days) HMXBs are highly circularised with eccentricity less than 0.006. But the orbit of 4U 1538–52 is not yet circularised and shows no orbital decay. Therefore, it is likely that 4U 1538–52 is a relatively young system compared to the other HMXBs. Using the mass, radius and luminosity of the companion star QV Nor from the optical observation (Reynolds *et al.* 1992) the approximate time for tidal circularisation of this system works out to be 2×10^3 yr (Lecar *et al.* 1976). Since the system is still eccentric, it is possible that the age of the system after the supernova explosion is of the order of a few thousand years; very small compared to the lifetime of HMXBs.

The eccentricity derived using the two RXTE observations in 1997 (Clark 2000) and 2003 (reported here) is likely to be correct for two reasons. The RXTE-PCA has a large effective area compared to the earlier missions and in the energy range of PCA more photons are detected compared to CGRO-BATSE using which Rubin *et al.* (1997) did not detect any eccentricity. If the eccentricity is high, then either the binary system is young or the circularisation time is actually much longer than a few thousand years estimated using Lecar's method (1976). On the other hand, a binary is unlikely to become very X-ray active by wind accretion within just a few thousand years of the supernova explosion. It is possible that the circularisation mechanism works very efficiently in systems with Roche-lobe overflow like LMC X-4 while accretion mechanism in 4U 1538–52 is likely to be driven by stellar wind.

Lack of any substantial variation in the photon index, the cut-off energy, and the e-folding energy over the binary phase of 4U 1538–52 suggests that the continuum X-ray spectrum of the pulsar is hardly affected during its revolution along the orbit. This indicates a nearly constant accretion rate and identical conditions in the accretion column throughout the binary orbit. In the HMXB pulsar GX 301-2, orbital phase resolved spectroscopy with RXTE-PCA showed that the spectral parameters remain almost constant throughout the orbit though the overall accretion rate and wind density, and clumpiness of the wind changes considerably with the binary phase (Mukherjee & Paul 2004). On the other hand, in 4U 1538–52 the equivalent hydrogen column density shows a smooth variation over the orbital phase. We compare the observed column density profile with a model estimated by assuming a spherically symmetric Castor *et al.* (1975) wind from the companion star. The velocity profile of the wind is:

$$v_{\text{wind}} = v_{\infty} \sqrt{1 - \frac{R_c}{r}} \quad (5)$$

where v_{∞} is the terminal velocity for the stellar wind, R_c is the radius of the companion and r is the radial distance from centre of the companion star. The column density profiles were derived using a numerical integration along the line of sight from the pulsar to the observer for three different inclination angles 65° , 75° and 85° respectively. In Fig. 7, we show the equivalent hydrogen column density measured with BeppoSAX and RXTE-PCA as a function of the orbital phase with respect to the mid-eclipse time. A mass-loss rate of $\sim 10^{-6} M_{\odot} \text{ yr}^{-1}$ and $v_{\infty} \sim 1000 \text{ km s}^{-1}$ were assumed. The model calculations of the absorption column density for different inclination angles when superposed on the observed values indicate reasonable agreement as can be seen in Fig. 7. This indicates that a spherically symmetric stellar wind from the companion

star may produce the observed orbital dependence of the column density for certain range of the orbital inclination.

We note here that we have not done any fitting of the column density measurements and the model column density values at different orbital phases (Fig. 7). For the wind density model used here, the phase resolved column density values depend on the mass loss rate, the terminal velocity and the inclination angle of the binary orbit. The data presented here is not suitable for such a detailed analysis but only shows that such a model is consistent with the observations. In eclipsing X-ray binaries like 4U 1538–52, the eclipse duration puts strong constraints on the orbital inclination. We point out that the orbital phase resolved column density measurements can be an independent way of estimating the orbital inclination, especially for non-eclipsing X-ray binaries. We also note that in several eclipsing and wind-fed HMXBs, the eclipse ingress is more gradual than the eclipse egress (Haberl *et al.* 1989; Feldmeier *et al.* 1996; Naik & Paul 2004). From the RXTE-ASM light curves of many eclipsing HMXB pulsars we have verified that this effect is more pronounced at lower energies. This ingress–egress asymmetry, which varies in degree from source to source is partly due to absorption by accretion material trailing the X-ray source in its orbit. For example, Vela X-1 (Feldmeier *et al.* 1996) has a much larger ingress–egress asymmetry and LMC X-4 (Naik & Paul 2004) has a much smaller asymmetry compared to 4U 1538–52 (Fig. 1). We think that the slight asymmetric accretion column distribution in the case of 4U 1538–52 on two sides of the eclipse (Fig. 7) can also partly be due to some trailing accreting material.

In several other HMXBs like 4U 1700–37 (Haberl *et al.* 1989), 4U 1907 + 09 (Roberts *et al.* 2001) and GX 301–2 (Leahy 1991; Pravdo & Ghosh 2001, with RXTE), it has been shown that a simple spherical wind emanating from the companion star is not sufficient to explain the column density profile. Density enhancement due to a trailing stream or a disk from the companion arising due to the dynamical effects of the pulsar was invoked to explain the said profile. In the case of GX 301–2, detailed observations with RXTE-PCA revealed that neither a disk nor a stream was sufficient to explain the observed profile (Mukherjee & Paul 2004). Considering the results for these HMXBs, the present work seems to carry an interesting implication. Clark *et al.* (1994) have studied the column density variation during eclipse egress for this particular source with GINGA data. Two HMXB pulsars in which similar increase in absorption column density pattern was seen near eclipse are X1908+075 (Levine *et al.* 2004) and SMC X-1 (Woo *et al.* 1995) which also may have isotropic wind pattern from the companion stars. In the present study we investigated the out-of-eclipse variation of the column density and found that a simple model with a spherically symmetric stellar wind describes the observations well. Since 4U 1538–52 has a moderate X-ray luminosity ($\sim 10^{36}$ erg s $^{-1}$), it is not expected to cause any significant perturbation in the acceleration of the stellar wind through X-ray ionization as compared to high luminosity sources like Cen X-3 (Day & Stevens 1993) and SMC X-1 (Woo *et al.* 1995) with luminosities of $\sim 10^{38}$ erg s $^{-1}$ in the high states. Observation of 4U 1538–52 covering the eclipse ingress and egress with high throughput telescope like XMM-Newton and SUZAKU which also have good low energy response will allow one to carry out a detailed study of the stellar wind structure in this source and also determine the extent to which the local X-ray photoionisation affects the stellar wind acceleration.

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Analysis of Enhanced Velocity Signals Observed during Solar Flares

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Abstract. Solar flares are known to release a large amount of energy. It is believed that the flares can excite velocity oscillations in active regions. We report here the changes in velocity signals in three active regions which have produced large X-class flares. The enhanced velocity signals appeared during the rise time of the GOES soft X-ray flux. These signals are located close to the vicinity of the hard X-ray source regions as observed with RHESSI. The power maps of the active region show enhancement in the frequency regime 5–6.5 mHz, while there is feeble or no enhancement of these signals in 2–4 mHz frequency band. High energy particles with sufficient momentum seem to be the cause for these observed enhanced velocity signals.

Key words. Sun: oscillations—Sun: sunspots—Sun: flares.

1. Introduction

Solar flares, which are associated with the active regions, release a large amount of energy within few minutes to hours. Apart from the high budget radiations, it also produces a large amount of energetic particles which move at high speeds. Since many years, several researchers have made efforts to understand the mechanism involved in these explosive events that occur in the solar atmosphere. The major aspect of such investigations involved looking for the change in magnetic flux (Lara *et al.* 2000), change in magnetic shear (Sakurai *et al.* 1992; Ambastha *et al.* 1993; Wang *et al.* 1994; Hagyard *et al.* 1999; Li *et al.* 2000), change in magnetic helicity (Moon *et al.* 2002) and change in magnetic energy (Metcalf *et al.* 2005), before and after the flares.

Apart from searching for the cause of the flare, people have also looked at the consequences of the flares in the solar atmosphere and one chief thrust area has been the study of the effect of flare on solar oscillations, especially on the *p*-mode oscillations frequency and power in the concerned active region. These global oscillations are present on the Sun all the time, peaking the power in 3 mHz frequency band in velocity oscillations. These global modes are excited by the turbulence in the convective layer. If this be the case on one side, researchers have found reduced velocity power in 3 mHz band in sunspots as compared to the quiet part of the Sun, whereas enhancement in high

frequency power in these active regions (Hindman & Brown 1998; Venkatakrishnan *et al.* 2002). Sunspots are known to absorb the p -mode power (Braun *et al.* 1987), however this situation may be different during the flares.

It is known that solar flares generally occur in active regions. Wolff (1972) suggested that large solar flares can stimulate the free modes of oscillation of the entire Sun. Haber *et al.* (1988) have reported that there is an average increase in the power of intermediate degree l , after the flare. At the same time, Braun & Duvall (1990) could not detect the excitation of acoustic waves from an X-class flare. Kosovichev & Zharkova (1998) have reported signature of seismic waves (solar quakes) during solar flare. Following this result, Donea *et al.* (1999) found an acoustic source associated with the flare using seismic images. They found a considerable flare induced acoustic power in 3.5 mHz band. The flare also induced power in 6 mHz band. Further, they report that 6 mHz flare signature lags the 3.5 mHz by ~ 4 min. Application of ring diagram analysis technique to several flare producing active regions showed that the power of p -modes appears to be larger when compared to that in non-flaring regions of similar magnetic field strength (Ambastha *et al.* 2003). Using helioseismic holography technique, Donea & Lindsey (2005) have reported emission of seismic waves from large solar flares which have occurred in active region NOAA 10486 on 2003 October 28 and 29. They found many compact acoustic sources associated with the foot-points of the coronal loops.

In this paper, we study the enhanced velocity signals in NOAA active regions 10486 and 9415 during the flare, co-temporal with the GOES soft X-ray flux. These active regions have produced major X-class flares and hence the effect of flare on velocity oscillations can be well studied. Also, each of the selected events are in decreasing order of flare strength. We have used Fast Fourier Transform based power spectral images to study the acoustic power in the active region in different frequency regimes during the flare event. The next few sections of the paper are devoted to the data and its analysis, while the later sections present a summary of the results and related discussions.

2. Data

In order to examine the enhanced velocity signals in active regions during the flare, we have used the full disk Dopplergrams observed at the cadence of 60 s, obtained from the Michelson Doppler Imager (MDI: Scherrer *et al.* 1995) on-board the Solar and Heliospheric Observatory (SOHO: Domingo *et al.* 1995). We chose the data of NOAA active regions 10486 and 9415 for the days 2003 October 28 and 29, and 2001 April 6, respectively, for our study. We have chosen these active regions because they produced large flares on the aforesaid days and hence it would be interesting to examine the effect of these flares on the acoustic oscillations. Complimented with these data sets, we also used the white light images and line-of-sight magnetograms from MDI, 1600 Å images from TRACE, soft X-ray flux data from GOES (1–8 Å) and hard X-ray flux data (100–300 keV) from RHESSI.

The active region NOAA 10486 was one of the largest groups of sunspots observed in the declining phase of solar cycle 23. It appeared on east limb of the Sun on 2003 October 23 in the southern hemisphere. This complex active region belonged to $\beta\gamma\delta$ and produced several X-class and M-class flares during its passage on the solar disk. The flare observed on 2003 November 4 was X28 flare, which was one of the largest flares ever observed till now from GOES X-ray detector, however during that time the active region was on the limb. The other flares were X17 and X10, which occurred on

Table 1. Active region numbers, date and time of observations, location of the active regions and the X-ray class of the flares are listed below.

Sl. no.	Active region no.	Date	Time (UT)	Location	X-ray class
1.	NOAA 10486	2003 Oct 28	06:00–14:59	S17E13	X17.2
2.	NOAA 10486	2003 Oct 29	16:00–23:59	S16W08	X10
3.	NOAA 9415	2001 Apr 06	16:20–23:59	S21E31	X5.6

2003 October 28 and 29, respectively, however during that time the active region was located near the central meridian.

The active region NOAA 9415 was compact sunspot group and belonged to $\beta\gamma$. This also appeared in the southern hemisphere and produced several X-class and M-class flares during its passage on the disk. The one we chose is X5.6 flare, which occurred on 2001 April 6 at 19:21 UT.

The NOAA no. of the active regions, date and time of observations, location of the active regions on the solar disk and the X-ray class of the flares are summarized in Table 1. The active region was tracked using heliographic co-ordinates in the full disk Dopplergrams and were interpolated to a $1''$ pixel size. The missing data sets were interpolated. We then used a two-point backward difference filter to remove the low frequency fluctuations. These filtered Dopplergrams were used for this study.

3. Analysis and results

3.1 Space–time diagrams

In order to identify the temporal changes in the velocity field fluctuations at different spatial locations, we used the space–time analysis technique. The space–time diagram for the time series of Dopplergrams has been obtained by selecting a pixel row along x -direction (as shown in Fig. 1(a)) and stacking it in time. To identify the location of the slit on the active region, we have shown the magnetogram and the position of the slit in Fig. 1(c). The magnetic field strength along the slit direction is shown for reference in Fig. 1(d). Figure 1(a) shows the space–time diagram for the 2003 October 28 event. This space–time diagram shows that there is a common mode of oscillations in the small magnetic field strength regions of the Sun. However, oscillations with reduced strength and of different kinds can be seen in the active parts of the Sun. The region with patchy strip is the location of the active region (between 80 and 260 arcsec) and has been confirmed from Fig. 1(c) and (d). In the space–time diagram (Fig. 1(a)), one can also notice that there is an apparent transient change in velocity signals (inside the box) in the active region during the flare (between 11:05 and 11:10 UT and between 100 and 250 arcsec). This can be clearly seen against the background oscillations as black and white patches from 120 to 180 arcsec, which is running horizontal to the x -axis in the magnified version of the space–time map (Fig. 1(b)). The GOES X-ray flux has been plotted adjacent to the space–time diagram to examine the time of transient changes in the velocity oscillations in the active region. These GOES X-ray data are recorded in the soft X-rays in the 1–8 Å band with a temporal resolution of 1 min and

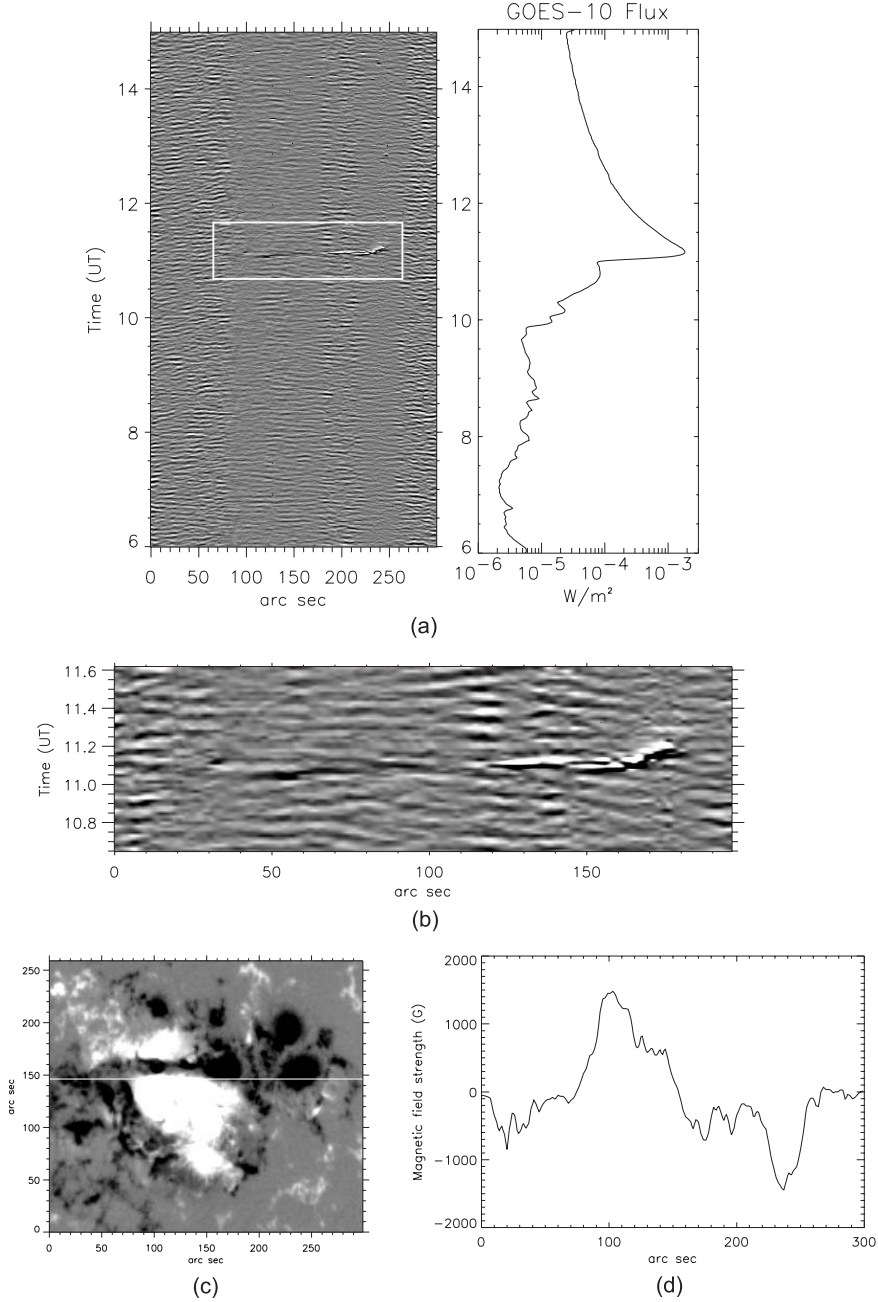


Figure 1. (a) Space-time map (left) showing the fluctuations in and around the AR 10486 obtained on 2003 October 28. The transient changes in velocity signals observed during the flare are shown inside the box. The GOES X-ray flux (right) is showing the peak time of the soft X-ray emissions. (b) A magnified version of the space-time map inside the aforesaid box to illustrate the transient velocity signals. (c) The magnetogram of the AR 10486 showing the position of the slit from which the space-time map has been made. (d) Magnetic field strength along the slit, showing the position of the strong and weak magnetic field regions.

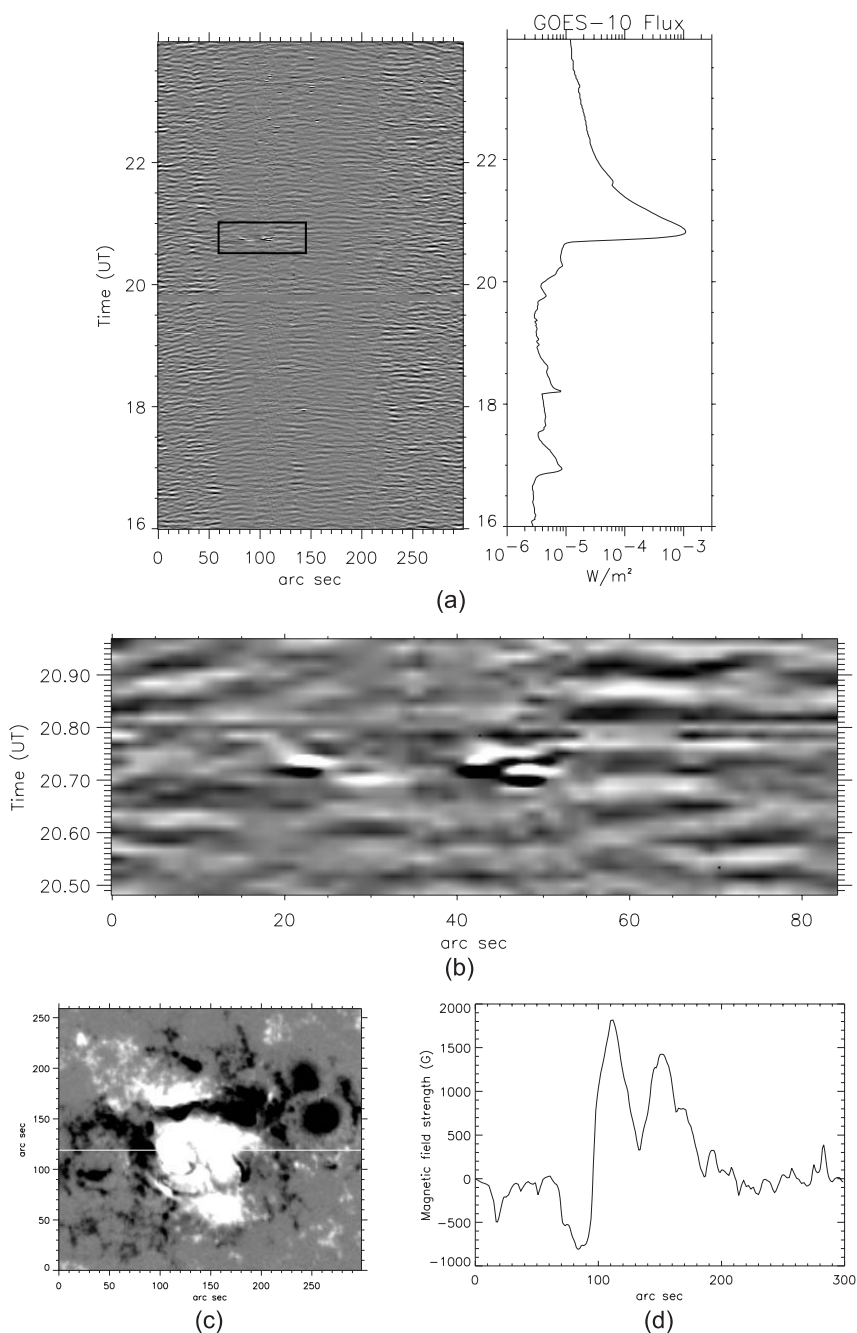


Figure 2. Same as Fig. 1, but for the flare event of 2003 October 29.

are good indicators of the magnitude of flares. The transient changes in the velocity signals appear to precede the peak time of the GOES X-ray flux.

Figure 2(a) shows the space-time diagram for 2003 October 29. The position of the slit is marked on the magnetogram (Fig. 2(c)). The boxed region in Fig. 2(a) shows

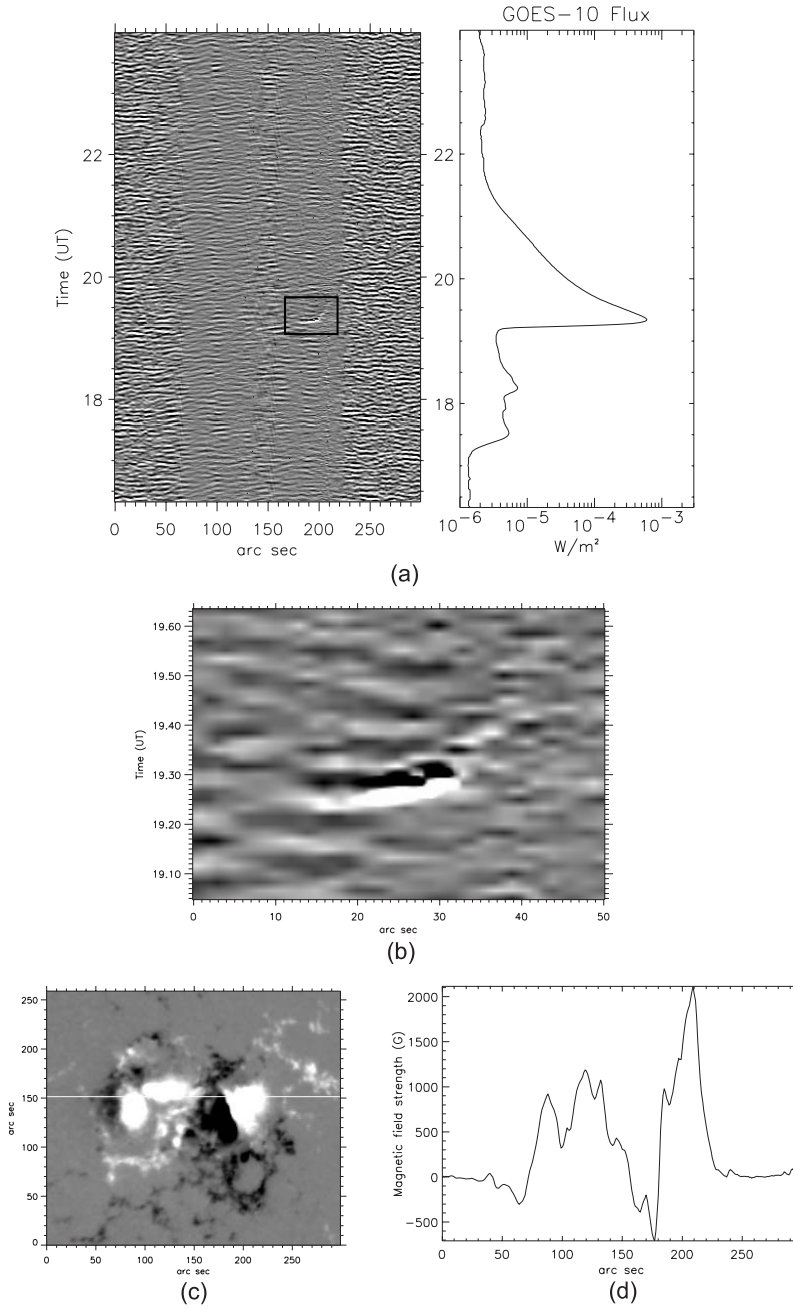


Figure 3. Same as Fig. 1, but for the flare event of 2001 April 6.

the transient change in the velocity oscillations and its magnified version is shown in Fig. 2(b). Figure 3(a) is the space-time map for 2001 April 6 flare event. Here also, in the magnified space-time map (Fig. 3(b)) one can notice enhancement in velocity signals in localized position of the sunspot, close to the peak time of the soft X-ray flux.

3.2 Power maps

In order to examine the enhancement of velocity signals in different frequency regimes and locations, we have computed the power maps of the active regions in selected frequency bands. It is well known that the typical distribution of acoustic power of photospheric oscillations shows a maximum at around 5 min (3.3 mHz). The band of dominant photospheric oscillations is in the regime of 2–4 mHz, with a decrease to negligible power at higher frequencies in the quiet Sun (Hindman & Brown 1998, and references therein). However, in active regions along with the reduced power in 2–4 mHz band, there is a little higher power in the high frequency regime (5–6.5 mHz) compared to the quiet part of the Sun (Venkatakrishnan *et al.* 2002, and references therein). In our study, we analyse the effect of flare on velocity oscillations in active regions. Therefore, we separate the aforesaid frequency bands for our analysis. The procedure for producing the power map is as follows: First, we took a data cube of filtered Doppler images of entire time series as mentioned in section 2. We then computed the temporal power spectra for individual pixels without averaging for time duration of two hours before the flare and similarly for two hours during and after the flare. At this stage, the time axis turns out to be the frequency axis. Now, we averaged the power images in the frequency regimes 2–4 mHz and 5–6.5 mHz, separately for pre-flare and post-flare sequences. The obtained power maps for different days of observation are shown in Figs. 4 and 5 for frequency regimes 2–4 mHz and 5–6.5 mHz, respectively. The left-hand side images show the average power map obtained before the flare while the right-hand side images are those obtained during and after the flare. In these power maps, the encircled regions are locations with enhanced power. The enhancement of power during the flare appears to be more pronounced in 5–6.5 mHz band as compared to 2–4 mHz band. In all these events, it is recognizable that there is little or no enhancement in the 2–4 mHz band. It is clearly seen that most of the locations in the active region appear dark in the power map, indicating absorption of power in the given frequency band. However, in the case of 2003 Oct 29 data, we do observe that some individual pixels in the quiet and active region appear bright before the flare in 5–6.5 mHz band. These could be noise appearing due to the bombardment of high energetic particles to the CCD pixels.

In order to estimate momentum in the enhanced power regions, we have used the integrated power over the pixels with enhanced signals (encircled regions shown in Fig. 5) in the 5–6.5 mHz band. For a typical density value in the line forming region $\approx 10^{-8} \text{ gm cm}^{-3}$, we have obtained momentum budget of $\sim 5.07 \times 10^{23}$, $\sim 4.32 \times 10^{21}$, and $\sim 1.22 \times 10^{22} \text{ dyn s}$ for 2003 October 28 and 29, and 2001 April 6 flare events, respectively, in the 5–6.5 mHz band. This value is close to the momentum budgets for smaller M-class flares as has been deduced by Zarro *et al.* (1988) and Canfield *et al.* (1990).

3.3 RMS map of velocity signals

In order to verify that the enhancement of the power obtained in power maps are due to the flare in origin, we computed the root mean square velocity map (rms map). These images were computed from the subtracted Dopplergrams. We have used the one hour data centered at flare occurrence time. First, the mean velocity image was obtained and then the rms image. In Fig. 6, the left-hand side images show the rms map of the

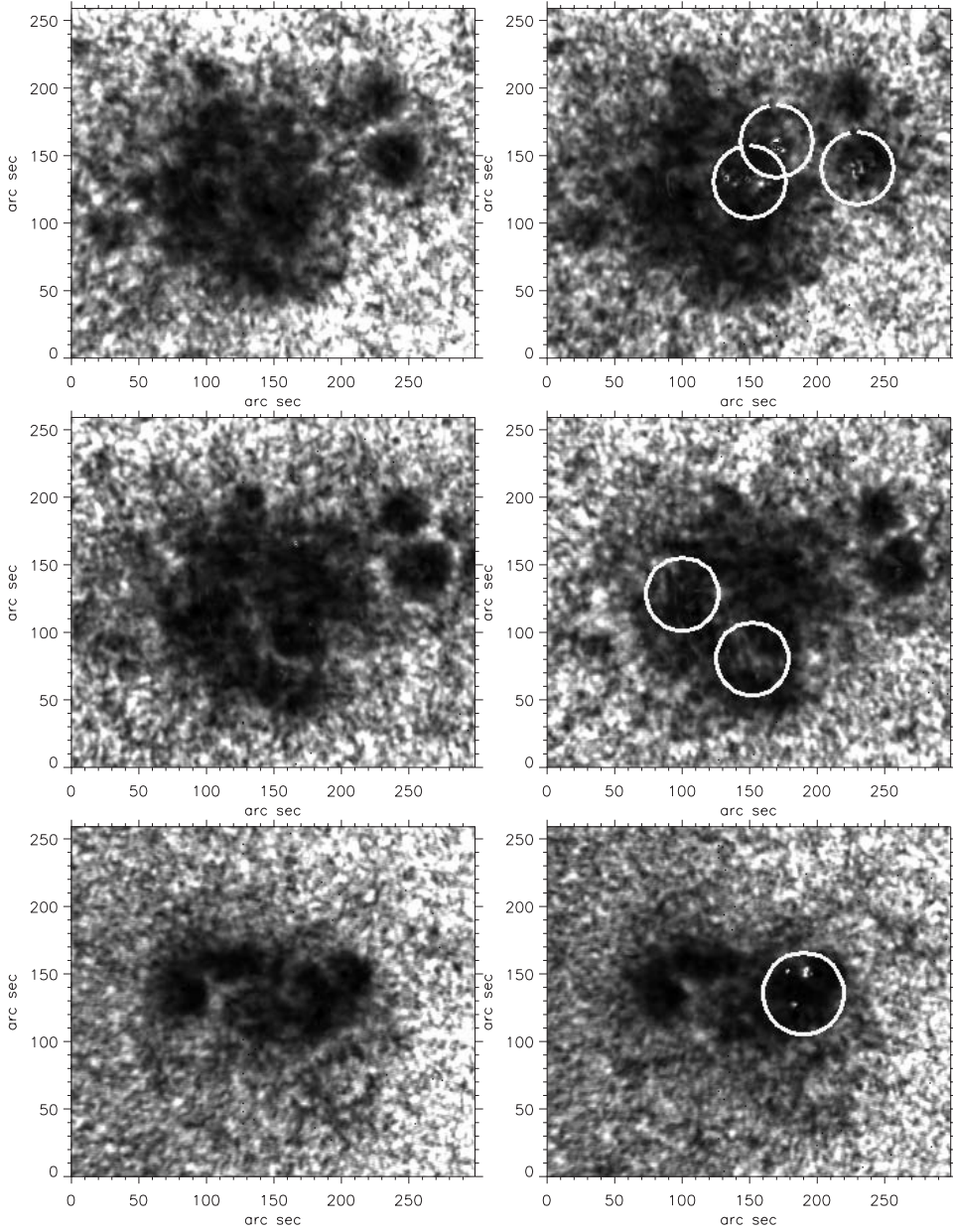


Figure 4. Power maps computed from the Dopplergrams for 2003 October 28 and 29, and 2001 April 6 events, from top to bottom, respectively. Left hand side images are power maps obtained by averaging the power in the frequency band 2–4 mHz for time duration of two hours before the flare, while the corresponding right hand side images are those obtained during and after the flare. The power maps on the left hand side have maximum (minimum) values of power 83.61×10^3 (129.6), 70.673×10^3 (419.65), and 85.371×10^3 (284.273) $\text{m}^2 \text{s}^{-2}$, while for the right hand side images the values are 81.05×10^3 (250.76), 69.834×10^3 (386.14) and 83.613×10^3 (405.88) $\text{m}^2 \text{s}^{-2}$, from top to bottom, respectively. The encircled regions show the location of enhanced power.

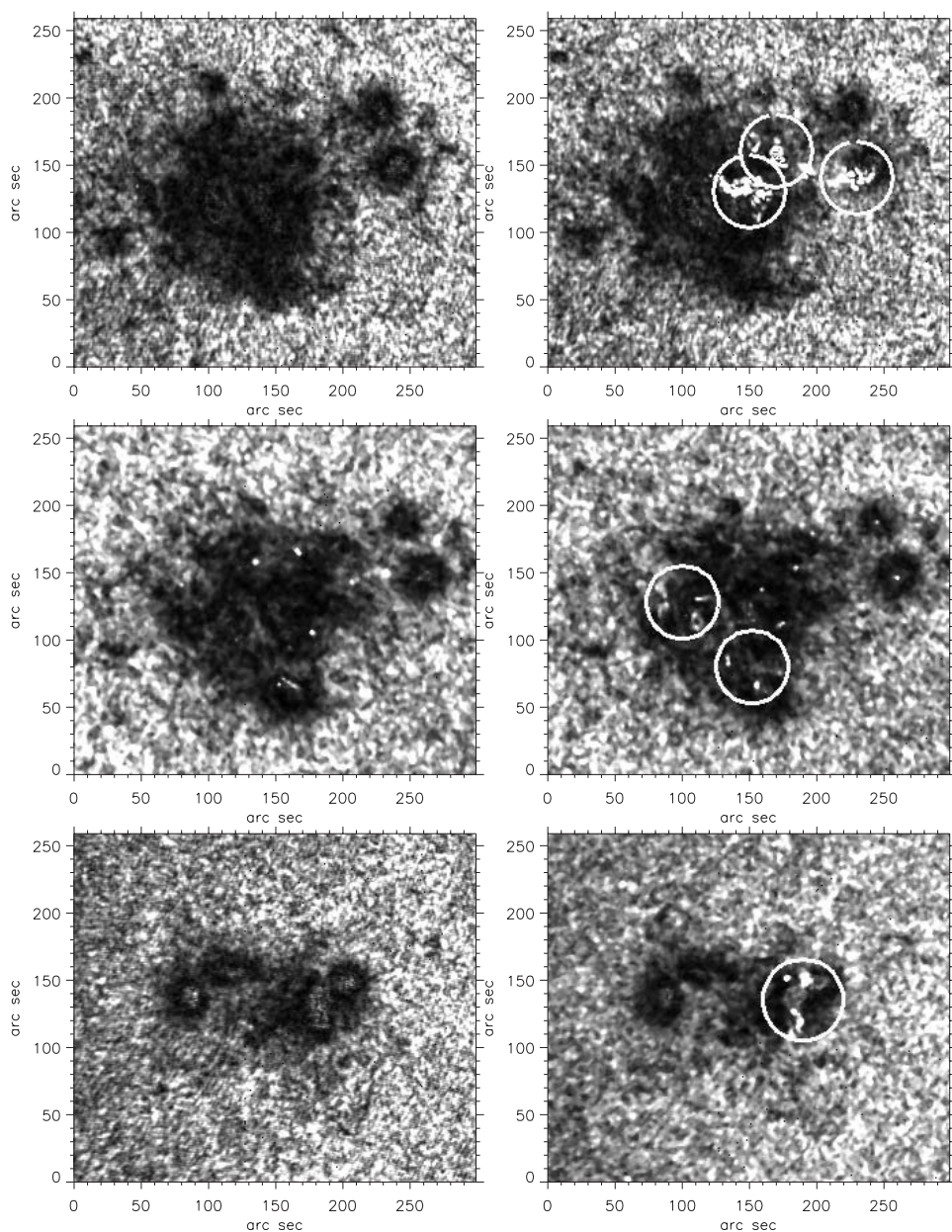


Figure 5. Same as Fig. 4, but for the frequency band 5–6.5 mHz. The power maps on the left hand side have maximum (minimum) values of power 25.54×10^3 (245.1), 114.185×10^3 (213.846), and 253.85×10^3 (242.896) $\text{m}^2 \text{s}^{-2}$, while for the right hand side images the values are 258.07×10^3 (199.358), 40.352×10^3 (275.154) and 182.748×10^3 (265.769) $\text{m}^2 \text{s}^{-2}$, from top to bottom, respectively. The encircled regions show the location of enhanced power.

velocity signals obtained for an hour before the flare and the right-hand side images are those for during the flare for the events of 2003 October 28 and 29, and 2001 April 6, from top to bottom, respectively. These rms maps show that there is enhancement of

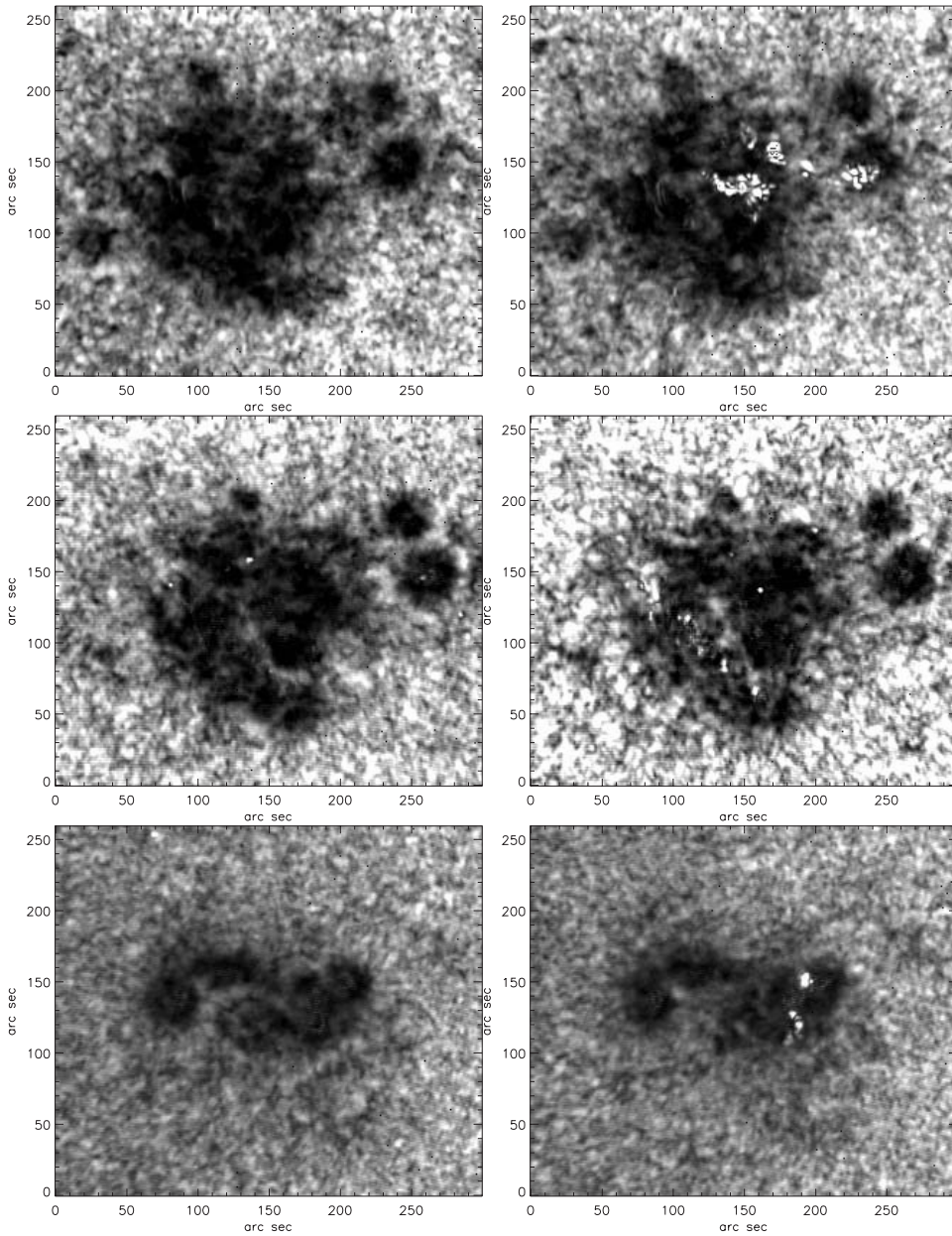


Figure 6. Root mean square (rms) images computed from the Dopplergrams for 2003 October 28 and 29, and 2001 April 6 events, from top to bottom, respectively. Left hand side images are the rms maps obtained for an hour before the flare events, while the corresponding right hand side images are obtained for an hour during the flare.

velocity signals at various locations of the active region during the flare events. This can be readily seen by comparing the rms maps obtained during the flare with those before the flare (left-hand side images). It should be noted that these enhancements

are exactly in the same position as that found in the power images of 5–6.5 mHz band. On 2003 October 28, these velocity enhancements look like two ribbons on either side of a neutral line, as usually seen in H_α observations of solar flares. These velocity enhancements are spatially located on one side between 130 and 160 arcsec in x -direction and between 120 and 140 arcsec in y -direction. The other enhanced region is located between 150 and 250 arcsec in the spatial x -direction and 130 and 160 arcsec in the spatial y -direction. Similarly, on 2003 October 29, the enhancement in the signal can be seen between 50 and 150 arcsec in the spatial x -direction, diagonal to the map. On 2001 April 6, the enhancement position is at 190 arcsec in the spatial x -direction. The advantage of this rms map is that it shows the changes which have occurred in time and also it shows the location of these changes in the active region. The rms map represents average power over the available acoustic frequencies. In the rms map, one can also notice that the sunspot region is dark compared to the surrounding region of the spot, which is because of the absorption of acoustic modes in the sunspots as compared to the quiet region. However, the locations of acoustic emissions in the active region are bright. The coincidence of the positions of the velocity enhancements in both the rms maps and power maps suggest that these enhancements in velocity signals are due to the flare.

In order to examine the positions of the enhanced velocity signals with respect to the hard X-ray positions, we used the RHESSI hard X-ray images in the range 100–300 keV. We then overlaid the contours of these hard X-ray signals on the rms images. Unlike the rms images obtained above, here we used the full-disk, derotated and subtracted Dopplergrams for rms image extraction. This is for easy alignment of the rms images with the RHESSI and TRACE images. We also overlaid the contours of hard X-ray on the TRACE 1600 Å images to examine the positions of the loops and the hard X-ray flux positions. Figure 7(a), (b) and (c) shows the contours of RHESSI hard X-ray flux on the rms velocity image, the TRACE 1600 Å image and MDI intensity image, respectively, for the 2003 October 28 flare event. Figure 7(d), (e) and (f) shows the same for the 2003 October 29 flare event. In the case of the 2001 April 6 event, the RHESSI images are not available and hence we overlaid the contours of the loops observed in TRACE 1600 Å on the rms image and MDI intensity image, which may be the position of the hard X-ray flux. Figure 7(a) and (d) shows that brightening in the rms image occurs near the foot points of the hard X-ray flux. Even in the case of the flare on 2001 April 6, the brightening in the rms image coincides with the loop brightening of TRACE 1600 Å image. On 2003 October 28, the enhanced velocity positions are not completely correlated with the hard X-ray positions. The reason behind the lack of good correlation could be explained as follows: the RHESSI started observing the flare (after its night) at 11:07 UT, however the flare had started much before this time (09:51 UT). Thus, RHESSI observations could have missed the positions of hard X-ray source regions during the peak time in the given energy band. The hard X-ray source region moves as the reconnection site moves (Krucker *et al.* 2003; Bogachev *et al.* 2005). Since the flare started much before the RHESSI observations whereas MDI covered the whole event at a cadence of 1 min, the different excited positions in velocity are observable in MDI data, however RHESSI might have missed the corresponding positions in the hard X-ray.

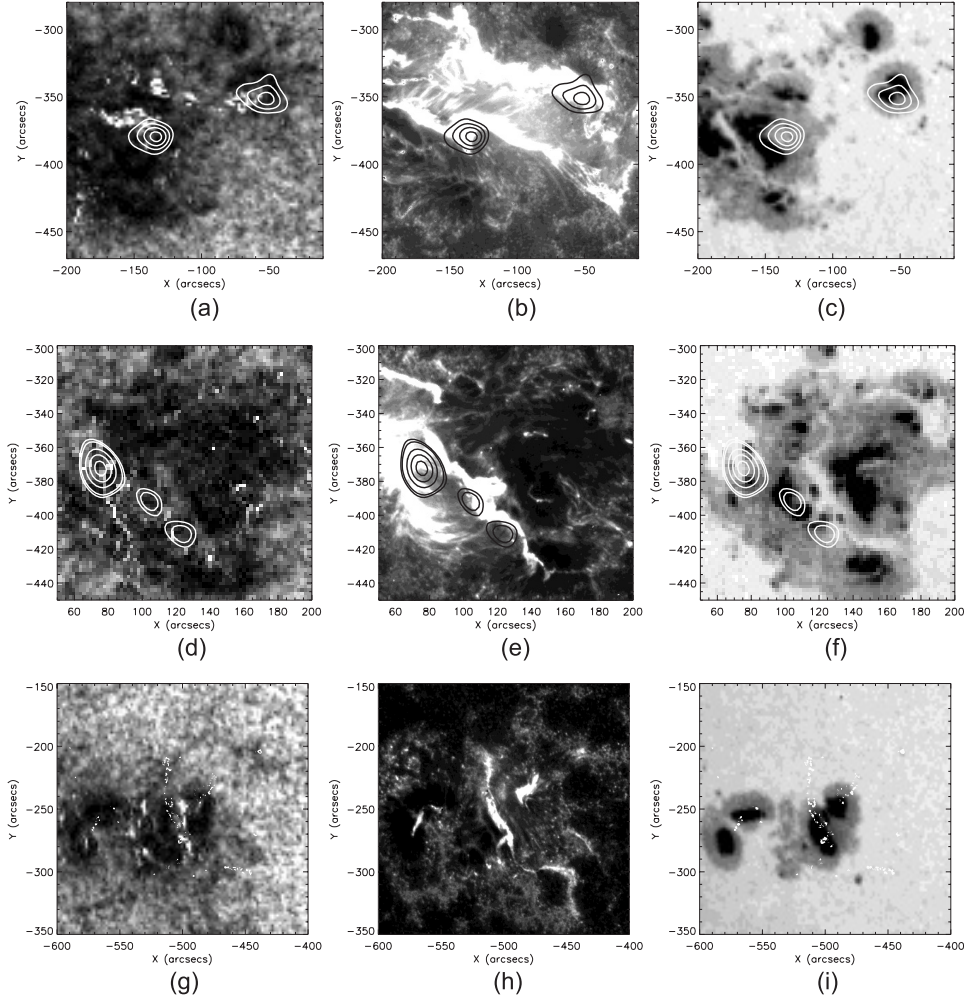


Figure 7. (a) The contours of RHESSI hard X-ray flux in the range of 100–300 keV, during the time interval 11:07–11:08 UT overlaid upon the rms image of 2003 October 28 flare. The contour levels are 90%, 70%, 50% and 30% of the maximum count. (b) The same contours overlaid upon the TRACE 1600 Å image observed at 11:28 UT. (c) RHESSI contours overlaid upon the MDI intensity image. (d), (e) and (f) are same as (a), (b) and (c) but for the data of 2003 October 29. The contour levels are 90%, 70%, 30% and 25% of the maximum count between the time interval 20:42–20:43 UT. The TRACE 1600 Å image was observed at 21:08:53 UT. (g) and (i) show the contours of the loop brightening observed in the TRACE 1600 Å flare overlaid upon the rms image and MDI intensity image of the 2001 April 6 flare event. (h) The TRACE 1600 Å image obtained at 20:03 UT is shown for comparison.

4. Summary and discussions

The study of velocity enhancements during X-class flares in the active regions NOAA 10486 and 9415 on 2003 October 28 and 29, and 2001 April 6, respectively, showed the following results:

- Transient change in velocity signals have been observed, which appear to precede the peak time of the GOES soft X-ray flux.
- These enhanced velocity signals are close to the vicinity of hard X-ray source regions.
- The enhancement in the velocity signal is found in the frequency regime 5–6.5 mHz, while little or no enhancement is observed in 2–4 mHz band.

The hard X-ray emissions have their origin from the non-thermal bremsstrahlung of accelerated electrons in solar flares (Brown 1971; Hudson 1972). The large flux of non-thermal energetic electrons can heat the chromospheric plasma (Fisher *et al.* 1985), and this sudden heating causes thermal instability which expands the chromospheric plasma to the temperature of the corona. This will happen when the hydrodynamic expansion time scale is larger than the heating time scale (Fisher *et al.* 1985). Further, the increase in plasma pressure will cause explosive evaporation of the chromospheric plasma into the corona and then downward settlement as chromospheric condensation (Fisher 1989), as has been evidenced in red wing asymmetry of H_α line profiles. The downward propagating plasma and shock waves ahead of plasma were assumed to be the agents delivering momentum to the photosphere which can result in observed excited acoustic signature. The excited velocity signals appear as red-shifted in those locations for short duration.

The global p -modes that are observed to be always present on the Sun are excited by turbulence in the convection zone (Goldreich & Kumar 1990, and references therein, Stein & Leibacher 1974) and are modified in active regions due to absorption of acoustic waves by magnetically inhomogeneous media (Braun *et al.* 1988). In addition to this continuous source of excitations, transient phenomena like flares can contribute additional energy to these modes. Even if a small fraction of the energy released by flare is transferred to oscillations in active region, it can increase the amplitude of oscillation significantly. However, the photosphere and magnetic field in active region provides a boundary condition for the p -modes and any source of additional energy from the external agent may affect these oscillations for short spans, as these cannot be sustained by the cavity. This can lead to the enhanced velocity oscillations during flare events.

On the other hand, there are some evidences about the Ni I 6768 Å line profile turning into emission during strong X-class flares (Qiu & Gary 2003). The non-thermal electrons can produce heating at the foot points of the magnetic field lines. Using non-LTE calculations, Ding *et al.* (2002) found that the Ni I 6768 Å line may appear in emission in the presence of strong electron beams in a cool atmosphere like sunspots. They also found that even if the atmosphere gets hot in the absence of energetic particles, the line profile may not turn into emission. If the line profile turns into emission during the flare, then there is a possibility of a sudden change in the velocity signal. This change in velocity signal persists till its source dies down. To examine this, one needs to have the spectral line information of line from which the Dopplergrams have been obtained. The future space-based missions like Solar-B may help in this context. In any case, one needs to study more about the velocity oscillations in active regions during the flares, since it may provide an exciting clue to the sub-photospheric flows as well as thermal anomalies in sunspots.

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Research Note

On New Limits of the Coefficient of Gravitation Shielding

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Abstract. New limits of the shielding coefficients in the supposed phenomenon of gravitation shielding have recently become available. The new values are briefly reviewed and discussed in order to update the state of art since some new limits for gravitation shielding are not necessarily the lowest ones which, instead, are those of interest when planning new experimental research or studying theoretically the possible effects of gravitation shielding.

Key words. Gravitation—shielding—lower limits.

In the theory of the supposed gravitation shielding Bottlinger (1912) suggested that a gravitational ray l of intensity g be weakened after crossing a layer of material with density δ according to the law

$$g = g_0 \exp\left(-\lambda \int_l \delta dl\right) \quad (1)$$

where λ is the so called shielding coefficient. The experiment made so far allows us to assume that $\lambda < 10^{-14} \text{ g}^{-1} \text{ cm}^2$ and for layer thickness and density sufficiently small, we may then write

$$g = g_0 \left(1 - \lambda \int_l \delta dl\right). \quad (2)$$

A straight-forward method to verify the reality of the phenomenon or to estimate λ is to measure the weakening of a gravitational ray caused by the crossing of a layer of known density and thickness. The first physicists who investigated this phenomenon and estimated λ , followed this path.

Many experiments have been made to verify the existence of gravitation shielding; a description of experiments made with a couple of horizontal pendulums and gravimeters may be found in Caputo (1977), articles with reviews may be found in the

more recent book edited by Edwards (2002). A review of conventional explanations of anomalous observations of presumed shielding made during solar eclipses is in the recent paper by Duif (2004).

To enhance the scope of this note we quote the experiments of Majorana (1919, 1920) who made a series of laboratory experiments performed with very refined techniques seeking to observe the variation of weight of a 1.3 kg lead sphere when it was screened from the effect of Earth's gravitational field with other masses. As screen he used a 114 kg mercury cylinder, then a 9.8 kg lead cube which completely surrounded the sphere.

From the results of these experiments Majorana (1919, 1920) was induced to state that the shielding effect existed. One experiment made with the mercury cylinder as shield led to conclude that $\lambda = 7 \cdot 10^{-12}$, while the experiments made with the lead cube concluded that $\lambda = 2 \cdot 10^{-12}$.

Following the publication of the results of Majorana's experiments, Russel (1921) showed that if we accept that the outer layers of celestial bodies shield the inner layers then the inertial mass of the planets would not be proportional to their gravitational mass and consequently, the motion of celestial bodies should differ notably from what is observed in reality. According to Russel (1921) this would condition the value of λ to be smaller than the values given by Majorana by a factor of 10^{-4} .

We consider here the three methods used for the detection of the gravitation absorption: using the Moon as a screen, using the Earth as a screen and observing the Moon motion fluctuations. Concerning the method in which the Moon is used as a screen it is of interest that the experimental limit set for λ by Slichter *et al.* (1965) with observations, taken with a La Coste tidal gravimeter during the solar eclipse of February 15, 1961 is $\lambda = 8.3 \cdot 10^{-16}$. During the same eclipse of February 15, 1961, from the data of the Great Horizontal pendulums of the Grotta Gigante in Trieste (Marussi 1960), Caputo (1962) set the limit $\lambda < 6 \cdot 10^{-16}$.

In a recent experiment, Yang & Wang (2002) observed possible gravity variation with a La Coste gravimeter during the total solar eclipse of March 9, 1997, and concluded that "no significant anomaly during the very solar eclipse was found. However, there are two gravity anomaly valleys with near symmetrical decrease of about $6-7 \mu\text{gal}$ at first contact and at last contact."

This interesting result was discussed by the authors in great detail and they concluded that the data set an upper limit for the existence of gravitation shielding. According to the formula (2) the anomaly recorded by Yang & Wang (2002), taking into account the elevation of the Sun at the time of the eclipse, would lead to an upper limit of λ to about $3 \cdot 10^{-14}$ (Yang & Wang 2002). We note that Yang & Wang (2002) used a gravimeter less sensitive by one order of magnitude than that used by Slichter *et al.* (1965) who obtained a lower limit for λ .

Results similar to that of Yang & Wang (2002) are obtained also from the data of Mishra & Rao (1997) who observed the solar eclipse of October 24, 1995 at Dhoraji with a gravimeter similar to that used by Slichter *et al.* (1965).

The method using the Earth as a screen was introduced by Harrison (1963) who used tidal gravimeters observation and found $\lambda < 10^{-15}$. A better result is that of Unnikrishnan *et al.* (2002) with $\lambda < 2 \cdot 10^{-17}$ who used the same method and analyzed 11,000 minutes of data taken by Wang *et al.* (2000) with a gravimeter during the 1997 total eclipse in China.

Finally studying the fluctuations of the Moon motion Crawley *et al.* (1974) found $\lambda < 6.3 \cdot 10^{-15}$. However, the most stringent limit on gravitation absorption, is that of Eckhardt (1990) who, in a brief note, using the Laser ranging to the Moon data of Williams *et al.* (1976), gave what is now the best upper limit $\lambda < 2 \cdot 10^{-21}$.

In conclusion, in the case of new research on gravitational shielding based on the use of the Moon or the Earth as screen and using gravimeters, considering the best results previously obtained for the coefficient of absorption and taking into account the elevation of the Sun, the instruments must be one order of magnitude more accurate than those previously used; moreover, considering the best result obtained by observing the anomalies in the Moon motion (Eckhardt 1990), we may not obtain better limits unless the gravimeters have an accuracy many orders of magnitude better than that of the instruments previously used.

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Interpretation of the Global Anisotropy in the Radio Polarizations of Cosmologically Distant Sources

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Abstract. We present a detailed statistical study of the observed anisotropy in radio polarizations from distant extragalactic objects. This anisotropy was earlier found by Birch (1982) and reconfirmed by Jain and Ralston (1999) in a larger data set. A very strong signal was seen after imposing the cut $|RM - \overline{RM}| > 6 \text{ rad/m}^2$, where RM is the rotation measure and $\overline{RM}$ its mean value. In this paper, we show that there are several indications that this anisotropy cannot be attributed to bias in the data. We also find that a generalized statistic shows a very strong signal in the entire data without imposing the RM dependent cut. Finally we argue that an anisotropic background pseudoscalar field can explain the observations.

Key words. Polarization—magnetic fields—galaxies: active—galaxies: high-redshift—cosmology: miscellaneous—elementary particles.

1. Introduction

Polarizations of radio waves from extragalactic sources undergo Faraday Rotation upon propagation through galactic magnetic fields. This effect provides very useful information about astrophysical magnetic fields (Zeldovich 1983; Vallée 1997). The amount of rotation is proportional to the magnetic field component parallel to the direction of propagation of the wave and to the square of the wavelength λ . The observed orientation $\theta(\lambda)$ of the linearly polarized component of the electromagnetic wave can therefore be written as

$$\theta(\lambda^2) = \chi + (RM) \lambda^2 \quad (1)$$

where the slope, called Faraday Rotation Measure (RM), depends linearly on the line integral of the parallel component of the magnetic field along the direction of propagation of the wave and χ is the intercept, also called the intrinsic position angle of polarization, IPA .

The observed polarization angle χ , after the effect of Faraday rotation is taken out of the data, is observed to be dominantly aligned perpendicular to the orientation axis of the galaxy (Gardner & Whiteoak 1966; Clarke *et al.* 1980). Let ψ denote

the orientation angle of the galaxy. Then the offset angle $\beta = \chi - \psi$ is found to be approximately equal to $\pi/2$ for most of the sources, i.e., the distribution of β over a large sample of sources is found to peak at $\beta \approx \pi/2$. Besides this dominant trend, a smaller peak is also found at $\beta \approx 0$. This suggests the existence of two populations, some with polarization position angles parallel and others perpendicular to the galaxy axis (Gardner & Whiteoak 1966; Clarke *et al* 1980).

In 1982, Birch empirically observed an angular anisotropy in the offset angle β , using a data set of 137 points. Birch's statistics were questioned by Phinney & Webster (1983) and it was pointed out that the significance of Birch's result can be significantly reduced if the experimental errors in β are taken into consideration. Phinney & Webster also suggested that the signal observed by Birch might result from the presence of bias in the data. Kendall & Young (1984) further investigated Birch's claim of cosmic anisotropy with more sophisticated statistics and using an updated version of Birch's data. They found that the statistics were not consistent with isotropy at 99.9% confidence level. Later Bietenholz & Kronberg (1984) repeated the calculations using single-number correlation test statistic, originally proposed by Jupp & Mardia (1980). This statistic also showed strong evidence of anisotropy in Birch's data with a confidence level of 99.98%. They went on to create an independent set of 277 points which, however, showed no signal of anisotropy. This led to a dismissal of Birch's results but left unresolved the puzzling fact that his data had contained a signal of anisotropy at such a high level of statistical significance.

The possible existence of anisotropy in radio polarizations was reanalysed by Jain & Ralston (1999). The authors collected an independent set of 361 points from all the available catalogues. This set included β values for 29 sources which were contained only in Birch's compilation. Jain and Ralston also considered the data set of 332 sources obtained after deleting these 29 objects. The authors found that both the data sets showed a statistically significant signal of anisotropy. The observed signal can be expressed as follows

$$2\beta - \pi \approx \Lambda \cdot R \quad (2)$$

where R is a unit vector in the direction of the source. The vector Λ represents the three parameters of this fit and points in the preferred direction of this dipole anisotropy. The unit vector $\hat{\Lambda}$ controls the distribution of $2\beta - \pi$ on the dome of the sky such that it is predominantly positive in the direction of the axis while it is predominantly negative in the opposite direction. As discussed later this axis points in the direction roughly opposite to the Virgo cluster. Since this is an axis, we can equally well consider it as pointing in the direction of the Virgo cluster.

Subsequently, several observations provide independent evidence for a global anisotropy. A large scale anisotropy has also been observed in the optical polarizations from distant quasars (Hutsemékers 1998; Hutsemékers & Lamy 2001; Jain *et al.* 2004). It was found that the optical polarizations are aligned with one another over very large distances of the order of Gpc. A very strong alignment was seen in a patch at large redshifts centered at the Virgo cluster (Hutsemékers 1998). Furthermore, it was found (Jain *et al.* 2004) that the polarizations of the large redshift, $z \geq 1$, data sample are aligned over the entire celestial sphere. The low l multipole of cosmic microwave background radiation (CMBR) also provides evidence for large scale anisotropy. As shown in de Oliveira *et al.* (2004), both the quadrupole and octopole show a pla-

nar structure, with the two planes being very closely aligned with one another. The perpendiculars to these planes are found to point approximately in the direction of the Virgo cluster. Ralston & Jain (2004) argued that all these observations present a very strong evidence in favour of a global anisotropy. They showed that there exists five different axes, namely the axes associated with the radio anisotropy, the optical alignment, the CMBR dipole, quadrupole and octopole, all pointing approximately in the same direction. The probability that any two of these axes show the observed level of coincidence by a random fluctuation was found to be less than 6%. The joint probability that all the 5 axes coincide by chance is extremely small and hence provides a very strong evidence in favour of a global anisotropy.

1.1 Likelihood analysis of radio anisotropy

Jain & Ralston (1999) tested the isotropy of radio polarizations as null hypothesis using Maximum Likelihood analysis by taking into account the transformation property of β . A typical null distribution for angular variables is the von Mises (vM) distribution

$$vM(\theta) = \text{const} \times e^{k \cos(\theta)} \quad (3)$$

where k is a constant and θ is the angular parameter which is taken here to be 2β . The factor 2 arises since β refers to the polarization angle and $\beta + \pi$ is identified with β (Ralston & Jain 1999). The joint distribution of the angle β and source positions was taken to be of the form

$$f(2\beta)C(2\beta, \mathbf{R})h(\mathbf{R}),$$

where $f(2\beta)$ is the marginal distribution of β which is assumed to be the von Mises, $h(\mathbf{R})$ is the angular distribution of the sources and the correlation *ansatz* is taken to be of the form

$$C(2\beta, \mathbf{R}) = \exp[\boldsymbol{\lambda} \cdot \mathbf{R} \sin 2\beta]. \quad (4)$$

Here $\boldsymbol{\lambda}$ represents the three parameters of the correlation *ansatz*. An alternate distribution, obtained by replacing

$$2\beta \rightarrow \Omega(\beta) = 2\beta + \nu \sin 2\beta \quad (5)$$

was also considered by Jain & Ralston (1999) since it was found to provide a better null fit to the data. The fit selected a negative value of the parameter $\nu \approx -1$ which leads to a more sharply peaked distribution in comparison to the von Mises. Such a sharply peaked function is also justified by the distribution of polarizations obtained under the assumption that the electric field components of the electromagnetic wave are normally distributed (Sarala & Jain 2001, 2002).

Likelihood analysis showed a strong signal of anisotropy using the transformation (5) in the joint distribution (4). The strongest signal was obtained after making the cut on the rotation measures so as to eliminate the data with RM in the range $|RM - \overline{RM}| \leq 6$, where $\overline{RM} = 6 \text{ rad/m}^2$ is the mean value of the rotation measure over the data sample. In Fig. 1, we show the histogram of the rotation measures for the 332 sources. It is found that the peak position as well as the mean value of RM is shifted from zero and lies approximately at 6 rad/m^2 . This is somewhat unexpected since for a large

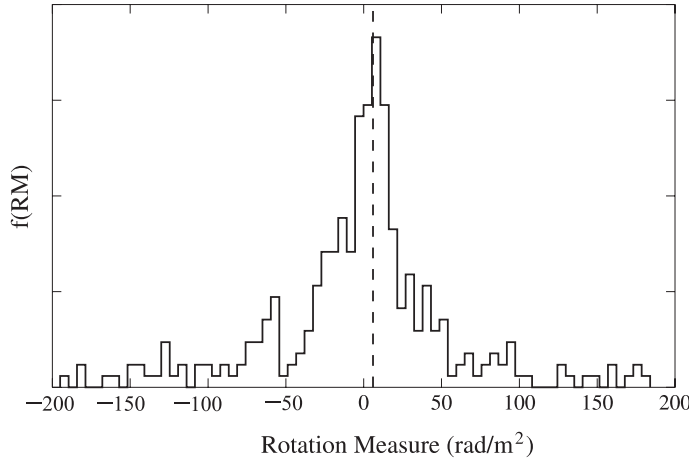


Figure 1. The distribution of Rotation Measure for the data with 332 points excluding some of the sources listed in Birch (1982) for which information on RM was not available. The RM distribution is not exactly Gaussian. Its mean as well as the peak position is shifted from $RM = 0$ towards positive values, as discussed in text. The vertical dashed line shows the position of the mean, $\bar{RM} = 6$.

unbiased sample we would expect the peak value of RM to be zero. Even for a larger data sample compiled in Broten *et al.* (1988) we find similar distribution with peak position as well as the mean and median of the rotation measures shifted from zero towards positive values. The cut on RM , therefore, removes points lying close to the central peak in the distribution. The statistical significance of the signal after this cut was found to be approximately 0.06%, i.e., the probability that the correlation seen in data might arise in a random sample is given by $P = 0.06\%$. The unit vector $\hat{\lambda}$ after making the cut $|RM - \bar{RM}| > 6$ was found to be (Jain & Ralston 1999) in the B1950 coordinate system:

$$\hat{\lambda} = [(0\ h, 9\ m) \pm (1\ h, 0\ m), -1^\circ \pm 15^\circ]. \quad (6)$$

Similar results were obtained by using the Jupp and Mardia correlation coefficient. In this case, the authors Jain & Ralston (1999) found the correlation between $\sin(2\beta - \sin 2\beta)$ and the angular position vector of the source. The parameter ν in Eq. (5) was set equal to -1 since the likelihood analysis generally preferred a value between -0.5 and -1 . The statistical significance in this case, after making the cut on rotation measures, was found to be $P = 0.04\%$.

In the present paper, we further illustrate the nature of this correlation. The basic cause of this correlation is so far unknown. There have been several proposals such as observational bias, bias in the extraction of RM and β (Phinney & Webster 1983) existence of a light pseudoscalar (Jain *et al.* 2002), cosmic rotation (Obukhov 2000; Obukhov *et al.* 1997; Mansouri & Nozari 1997), inhomogeneous and anisotropic universe (Moffat 1997), weak lensing by large-scale density inhomogeneities (Surpi & Harari 1999), background torsion field (Dobado & Maroto 1997) etc. The fact that the correlation is improved significantly after making a cut based on rotation measure may provide a hint about its origin. One obvious possibility is that there may exist

Table 1. JM correlation between RM and β with different cuts in the RM .

Cuts	n	$n\rho^2$
Full data	332	1.742
$ RM - 6 > 6$	265	1.780
$ RM < 100$	278	4.341
$ RM < 70$	254	1.101
$ RM < 100$ and $ RM - 6 > 6$	211	4.320

bias in the extraction of RM and β which leads to a correlation between these two parameters. Then, since the RM or the magnetic field is not isotropically distributed, it can induce an anisotropy in β . Indeed such a possibility was suggested (Phinney & Webster 1983), soon after the appearance of Birch's paper.

2. Statistical analysis of the angular correlation in polarizations

We next perform several statistical tests in order to determine whether the observed correlation can be explained in terms of bias in the values of β which may result from biased extraction of rotation measures. As discussed above, if the extraction of RM and β is biased then since the RM is not isotropically distributed on the celestial sphere it can lead to an anisotropic distribution of β . In order for this mechanism to work it is necessary that β is strongly correlated with RM . We next examine if such a correlation exists in the data by computing the Jupp & Mardia (1980) (JM) correlation between RM and β for the data set containing 332 sources. The correlation is found to be very small with $n\rho^2 = 1.742$. We tried several cuts to see whether the correlation is enhanced and found a small correlation with $n\rho^2 = 4.341$ for $|RM| < 100 \text{ rad/m}^2$ with the number of sources $n = 278$. A more stringent cut on RM does not improve the correlation. We summarize JM correlation with different cuts in Table 1. We see from this table that RM is not correlated with β . This implies that a large range of explanations of the anisotropy which invoke the possibility of biased estimation of RM and IPA , including the mechanism proposed by Phinney & Webster (1983), are disfavoured.

The above lack of correlation between RM and β , however, does not rule out the possibility that the value of β is randomized due to biased extraction of RM . This might happen preferentially in the regions of large magnetic field, i.e., where RM is large. As we have discussed earlier, the distribution of β peaks at $\pi/2$. If in a certain patch the values of β are randomized, then the mean value of β in this patch will deviate significantly from $\pi/2$ and could in principle give rise to a large value of the statistic indicating correlation.

We next investigate whether the correlation arises due to such a randomization. In Fig. 2, we plot the sign of $\sin 2\beta$ over the dome of the sky in the equatorial coordinate system after making the cut $|RM - 6| > 6$. Here the sources with $\sin 2\beta < 0.0$ and $\sin 2\beta \geq 0.0$ are represented by plusses and crosses respectively. From this figure we see that the positive values of $\sin 2\beta$ dominate near the dipole anisotropy axis given

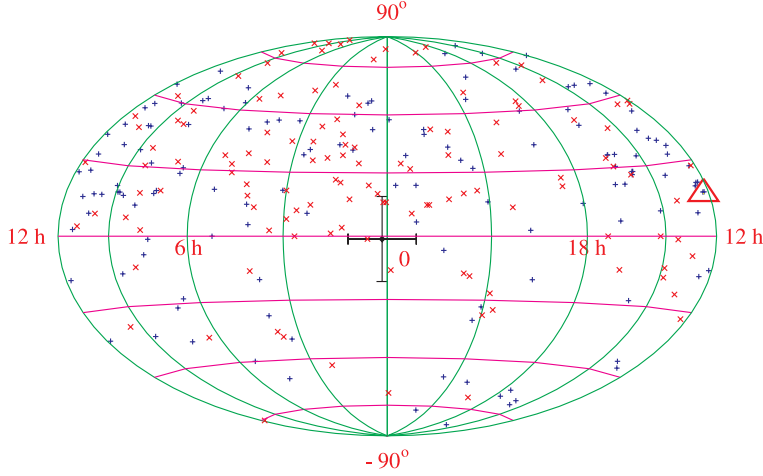


Figure 2. Aithoff–Hammer equal-area plot of $\sin 2\beta$, where β is the difference in the polarization position angle and the orientation angle of the source. We represent $\sin 2\beta < 0.0$ by blue pluses while $\sin 2\beta \geq 0.0$ is represented by red crosses. The dipole anisotropy axis, $[(0\ h, 9\ m) \pm (1\ h, 0\ m), -1^\circ \pm 15^\circ]$, obtained by Jain & Ralston (1999) is shown along with its error bars. The position of the Virgo supercluster is shown with a red triangle. The cut $|RM - 6| > 6$ has been used in this plot.

in equation (6). Figure 2 clearly indicates that there indeed exists an anisotropy in the polarizations of radio galaxies and there exist large patches where the value of $\sin 2\beta$ is either positive or negative. Hence we find no evidence that the correlation seen in data arises due to randomization of β in some regions. We may quantize this clustering of negative and positive values of $\sin 2\beta$ by defining a new statistic. We consider an N_n number of nearest neighbours of the source i . For these N_n sources we set $x_i = 1$ if $\sin 2\beta \geq 0$ and $x_i = -1$ if $\sin 2\beta < 0$. We then compute $X_i = |\sum_i x_i|$ and the statistic X is obtained by summing X_i over the entire data sample. The P -value in this case is computed by comparing with a large number of random samples obtained by shuffling the β values of the data. The resulting P -values or the significance level for the complete data sample and for the set obtained after making the cut $|RM - 6| > 6$ are shown in Fig. 3. We find that $P \approx 0.2 - 0.5\%$ for the complete sample and $P \approx 2.5 \times 10^{-4} - 6.0 \times 10^{-4}\%$ after the cut. In Fig. 4, we show the histogram of the statistic X generated using 10,000 random data samples. This figure also clearly shows that the statistic of the data sample lies far above the peak value of the histogram. The very small P -values obtained after the cut is a clear indication that β values in the different regions of the sky are indeed correlated with one another. This is infact another confirmation of the large correlation found by Jain & Ralston (1999). Hence we dismiss the above explanation in terms of bias in the extraction of IPA .

An improved statistical procedure for extraction of rotation measures and IPA was proposed by Sarala & Jain (2001, 2002). Using the revised β values obtained by this procedure, the Jupp–Mardia test for anisotropy gives $P = 0.1\%$. Hence these revised β values continue to support the presence of anisotropy in the data.

In order to further examine the effect of rotation measures on the observed anisotropy we examine how the correlation changes if we eliminate the sources which lie close

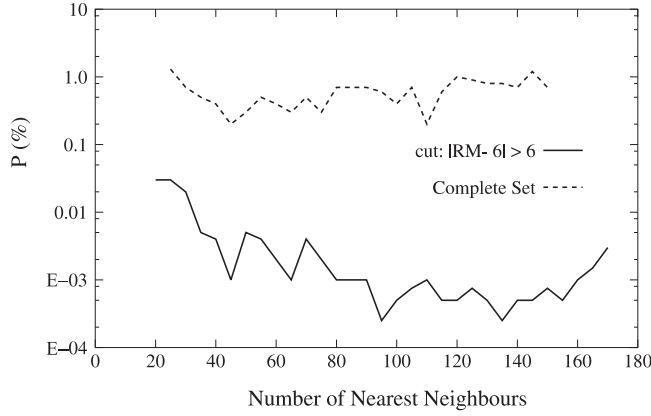


Figure 3. The P -value(%) using the X statistic is plotted as the function of the number of nearest neighbours. The result is shown for both the complete set (332 points) and with the cut in the RM (265 points).

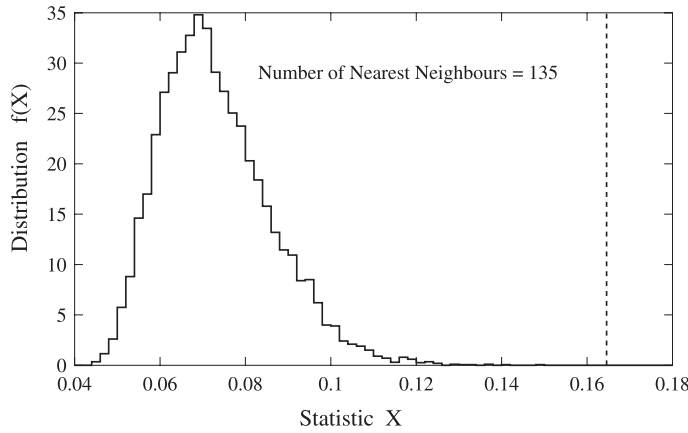


Figure 4. The histogram of X statistic using 10,000 random samples obtained by shuffling the β values of the data with the cut $|RM - 6| > 6$. The number of nearest neighbours is taken to be 135 at which the P -value is found to be the smallest. The statistic of the data in this case is equal to 0.165, shown by the dotted line, and the corresponding $P = 2.5 \times 10^{-4}\%$.

to the galactic plane. The dominant contribution to RM is obtained from the milky way and hence the largest values of RM are observed for these sources. We therefore impose a cut on the data to eliminate sources with the galactic latitude $|b| \leq 30^\circ$. This leaves a total of 214 sources in the data set. We repeat the JM correlation discussed earlier using Map 3. We find that the statistic $n\rho^2 = 9.43$ after the galactic cut and this implies that the anisotropy in β persists with about 98% confidence level. As expected, the anisotropy is enhanced by making a cut in the RM such that only the sources which satisfy $|RM - 6| > 6.0$ are included. The results are tabulated in Table 2. We see from the table that the statistic does not change much in comparison with the data without the galactic cut. Hence we find that the regions of large RM do not necessarily imply large correlation of β .

Table 2. The $n\rho^2$ values after eliminating sources lying in the galactic plane. Results are given for the full set as well as with a cut in the Rotation Measure.

	n	$n\rho^2$	$P(\%)$
Full set	214	9.43	2
$ (RM - 6) > 6.0$	155	22.33	0.04*

*The P value for the cut $|(RM - 6)| > 6.0$ is obtained by using χ^2_5 .

The above discussion clearly shows that the correlation seen in the data is not a direct consequence of the correlation of the RM with the angular coordinates of the source. However the cut in RM does result in a large increase in the signal of anisotropy and hence in some sense it must depend on RM . One possibility is that the sources eliminated by the cut, show a random behaviour and are uncorrelated. This turns out to be not true. Instead we find that the sources eliminated by the cut, show an opposite angular dependence in comparison to the remaining sources. We illustrate this by making the following *ansatz* for the joint distribution

$$F(\beta, \mathbf{R})h(\mathbf{R}) = \exp[k \cos \Omega + \lambda \cdot \mathbf{R}(1 + g(RM)) \sin \Omega] h(\mathbf{R}) \quad (7)$$

where as defined earlier $\Omega = 2\beta + \nu \sin 2\beta$ and $g(x)$ is taken to be a gaussian,

$$g(x) = c_1 \pi \exp \left[\frac{-(x - c_2)^2}{c_3^2} \right]. \quad (8)$$

Here we have introduced three new parameters c_1 , c_2 and c_3 . The difference of maximum likelihood between this correlated *ansatz* and the null *ansatz* turns out to be 17.6 for the entire data set of 332 sources. Here the null *ansatz* is the two-parameter distribution obtained by using the transformation (5) in the von Mises distribution. This “sharply peaked” distribution can be written as

$$S(\beta) = \text{const} \times e^{k \cos(2\beta + \nu \sin 2\beta)}. \quad (9)$$

The statistical significance of the signal of correlation in this case is

$$P = 3.8 \times 10^{-4}\%,$$

taking into account the eight parameters in the correlation *ansatz*. If the null hypothesis is assumed to be the von Mises distribution, we find an even smaller value of P . The signal of correlation is clearly very strong and cannot be dismissed as a statistical fluctuation. The parameters are given by $k = -0.561 \pm 0.071$, $\nu = -1.09 \pm 0.17$, $c_1 = -1.71 \pm 0.32$, $c_2 = 9.64 \pm 0.47$, $c_3 = 2.89 \pm 0.55$, $|\lambda| = 0.70 \pm 0.15$ and the axis parameters are given by

$$RA = (0^{\text{h}}, 3.7^{\text{m}}) \pm (0^{\text{h}}, 46^{\text{m}}), \quad DEC = -10.7^\circ \pm 12.2^\circ \quad (10)$$

in the B1950 coordinate system. Within errors this points opposite to the direction of the Virgo supercluster. We see from the value of the parameter c_1 that the sources whose RM lies roughly within the interval $c_2 \pm c_3$ show a correlation with $\lambda \cdot \mathbf{R}$ which is opposite to the remaining sources. The correlation term $\lambda \cdot \mathbf{R}(1 + g(RM))$ has an opposite sign for these sources in comparison to the remaining sources. The axis parameters, equation (10), obtained from this fit agree with the earlier fit, equation (6), within errors.

The function $g(RM)$ essentially treats the sources whose RM lies in the interval $c_2 \pm c_3$ differently from the remaining sources. This interval is centered at the value c_2 which is very close to the position of the peak in the RM distribution as can be seen in Fig. 1. We may attribute this shift in the peak to some bias in the extraction of RM or due to some systematic physical effect such as the contribution due to the milky way. Hence, if the difference $|RM - c_2|$ is small it may indicate that the host contribution and the contribution due to intergalactic propagation to RM is negligible. The introduction of the function $g(RM)$ in the correlation *ansatz*, therefore, may be interpreted as a dependence of the anisotropy effect on the host and the intergalactic magnetic fields.

The relationship between β and the position vector $\mathbf{R}$ of the source indicated by equation (7) may be expressed as follows,

$$\frac{\langle (1 + \nu \cos 2\beta) \sin \Omega \rangle}{\langle (1 + \nu \cos 2\beta) \cos \Omega \rangle} = \frac{\lambda \cdot \mathbf{R}}{k} (1 + g(RM)) \quad (11)$$

where the angular brackets indicate statistical averages. Since the parameter $\nu \approx -1$, the function $(1 + \nu \cos 2\beta)$ is approximately 2 near the peak of the distribution of β and zero near the tail. Hence this function weights the peak region more than the tail in the statistical averages. For β close to its peak value $\pi/2$, we find that $\sin \Omega \approx -2(1 - \nu)(\beta - \pi/2)$ and $\cos \Omega \approx -1$. If RM is not too close to its peak value or more precisely $|RM - c_2| > c_3$ then $g(RM) \approx 0$ and hence we find the approximate relation

$$\langle \beta - \pi/2 \rangle \approx \frac{1}{2(1 - \nu)} \frac{\lambda \cdot \mathbf{R}}{k}, \quad (12)$$

which is same as the relationship given in equation (2) with the Λ in equation (2) identified with $\frac{\lambda}{(1 - \nu)k}$ in equation (12).

The anisotropic distribution of the offset angle obtained with the fit equation (7) may now be illustrated by making an Aithoff–Hammer scatter plot of the correlation variable $(1 + g(RM)) \sin \Omega$. This is shown in Fig. 5 where stars and plusses represent sources for which $(1 + g(RM)) \sin \Omega > 0.1$ and $(1 + g(RM)) \sin \Omega < -0.1$ respectively. We find that the stars are concentrated in the direction of the axis with the plusses concentrated in the opposite direction. The anisotropy in the angular distribution of the offset angles is clear from this plot. The nature of the anisotropy is also nicely illustrated by making a scatter plot after taking the local averages. In Fig. 6, we show the Aithoff–Hammer scatter plot of the averaged correlation variable, $\overline{(1 + g(RM)) \sin \Omega}$, at the location of each source. The average is done over the nearest neighbours whose angular separation from the source is less than $\cos^{-1} 0.95 \approx 18^\circ$. The mean number of sources which contribute to this average is found to be 12.6. The plot clearly shows anisotropic nature of the distribution of the offset angles.

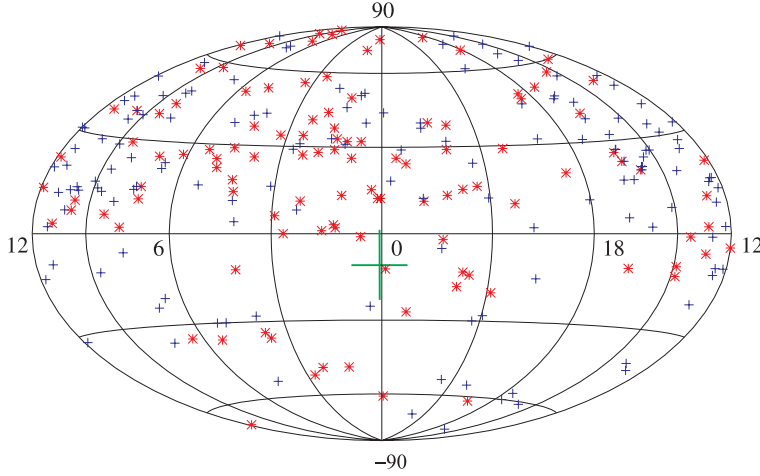


Figure 5. Aithoff–Hammer scatter plot of the correlation variable $(1 + g(RM)) \sin \Omega$, as defined in equation (7). Here red stars and blue pluses represent sources for which $(1 + g(RM)) \sin \Omega > 0.1$ and $(1 + g(RM)) \sin \Omega < -0.1$ respectively. The anisotropy axis, $[(0^h, 3.70^m) \pm (0^h, 46^m), -10.7^\circ \pm 12.2^\circ]$, along with its error bars is shown in green.

3. Physical origin of the effect

From the results obtained above we see that the offset angle β shows some correlation with RM but the correlation is direction-dependent. It is difficult to visualize how such a direction-dependent correlation can arise due to bias in data. In all likelihood this is caused by some physical phenomenon. The offset angles of the different sources may be intrinsically correlated or the effect may arise due to propagation. The possibility of intrinsic correlation can be further investigated by determining the correlation of the intrinsic position angle of polarization χ and the orientation angle of the galaxy ψ among the different sources. By using the statistical techniques used in Hutsemékers (1998); Hutsemékers & Lamy (2001); and Jain *et al.* (2004) to test the alignment of optical polarizations, we find that χ and ψ for different sources show no alignment with one another. The P value in this case is found to be larger than 10% independent of the number of nearest neighbours used to test for alignment. Hence it is not possible to attribute the radio anisotropy to intrinsic alignment of radio polarization and it is likely to arise due to some propagation effect. Within conventional physics the corrections to the Faraday polarization rotation effect are very small and hence it is unlikely that this anisotropy may be explained within this framework. The anisotropy may, therefore, be an indication of some new physical effect.

One interesting possibility is the presence of a very low mass background pseudo-scalar field in the universe. We assume that at large redshifts, relevant for the sources considered in the present data set, this field has a coherent component which is approximately given as

$$\phi(\mathbf{R}) \approx \phi_0 \cos(\theta)(1 + g[B_r(\mathbf{R})]) \quad (13)$$

where ϕ_0 is a constant independent of redshift, θ is the angular position of the source measured with respect to the preferred axis λ , B_r is the radial component of the

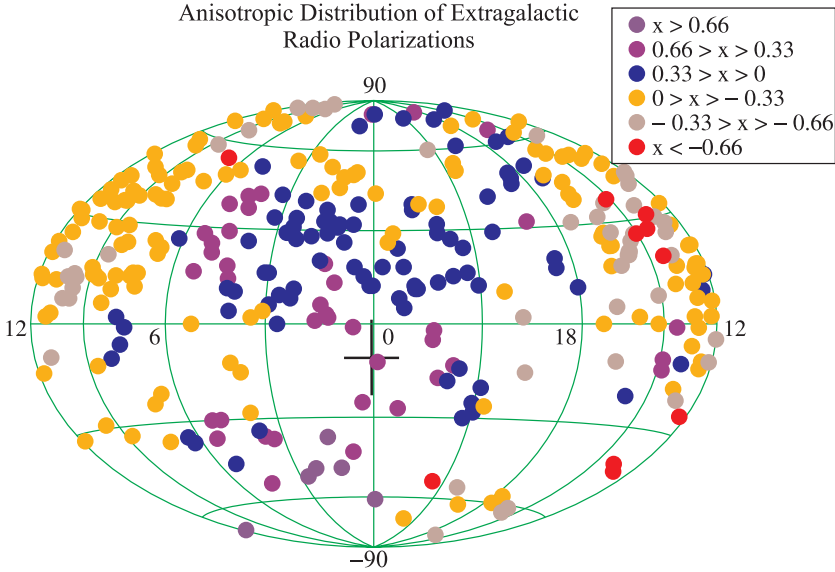


Figure 6. Aithoff–Hammer scatter plot of the averaged correlation variable $x = (1 + g(RM)) \sin \Omega$, as defined in equation (7). For each source located in the direction $\mathbf{R}_1$ we take the average over all the nearest neighbours which lie within the cone defined by $\mathbf{R}_1 \cdot \mathbf{R}_2 < 0.95$, where $\mathbf{R}_2$ is the unit vector in the direction of any other source. The anisotropy axis, $[(0^h, 3.70^m) \pm (0^h, 46^m), -10.7^\circ \pm 12.2^\circ]$, is shown along with its error bars. Within errors it points opposite to the Virgo supercluster.

magnetic field in the coordinate frame with origin located at the observation point, i.e., Earth and $g(B_r)$ is a gaussian defined in equation (8). At the position of the Earth we assume that this field acquires a value ϕ_1 . The total rotation in the polarization angle due to this field is then equal to $g_\phi(\phi(\mathbf{R}) - \phi_1)$, where g_ϕ is the coupling of pseudoscalar and electromagnetic fields (Harari & Sikivie 1992). Since the rotation measure is proportional to B_r we find that this effect can explain the observed correlation given in equation (7). This mechanism requires a large scale anisotropy in the universe and provides the simplest explanation of the observations.

The optical alignment effect may have some relationship to the radio anisotropy discussed in this paper since both the effects appear to be very strong in the direction of the Virgo supercluster. This effect is also not easily explained within conventional cosmology/astrophysics. As shown in Jain *et al.* (2002) the optical alignment effect can also be explained if we assume the existence of a light pseudoscalar. The proposed explanation requires the existence of several patches of large scale magnetic fields of length scales of order Gpc. Furthermore, the global alignment at large redshifts requires the existence of a magnetic field coherent over the entire universe. Finally the alignment of low l CMBR multipoles may also have some relationship with radio anisotropy.

4. Conclusions

To conclude, we find that there is considerable evidence of the presence of angular correlation in the radio offset angles which is not easily explained within conventional

cosmology/astrophysics. The observations may be explained by the presence of a hypothetical pseudoscalar particle. This effect, along with several other independent observations, indicate the existence of a global anisotropy in the universe with the preferred axis pointing approximately in the direction of the Virgo supercluster.

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Long-term Modulation of Cosmic Ray Intensity in relation to Sunspot Numbers and Tilt Angle

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Abstract. A detailed correlative analysis between sunspot numbers (SSN) and tilt angle (TA) with cosmic ray intensity (CRI) in the neutron monitor energy range has been performed for the solar cycles 21, 22 and 23. It is found that solar activity parameters (SSN and TA) are highly (positive) correlated with each other and have inverse correlation with cosmic ray intensity (CRI). The ‘running cross correlation coefficient’ between cosmic ray intensity and tilt angle has also been calculated and it is found that the correlation is positive during the maxima of odd cycles 21 and 23. Moreover, the time lag analysis between CRI and SSN, and between CRI and TA has also been performed and is supported by hysteresis curves, which are wide for odd cycles and narrow for even cycles.

Key words. Cosmic ray intensity—sunspot numbers—tilt angle.

1. Introduction

The long-term variations of galactic cosmic rays have been compared with the behaviour of different solar activity indices and heliospheric parameters several times (Belov *et al.* 2002 and references therein). The magnetic field on the source surface determines the structure and properties of the heliomagnetosphere. Therefore, it should be more closely connected with cosmic ray modulation than with other solar characteristics (sunspot numbers or coronal emission intensity). The intensity of galactic cosmic rays varies inversely with sunspot numbers, having their maximum intensity at the minimum of the 11-year sunspot cycle (Forbush 1954, 1958). The cosmic ray intensity curve also appears to follow a 22-year cycle with alternate maxima being flat-topped and peaked as predicted by models of cosmic ray modulation based on the observed reversal of the Sun’s magnetic field polarity after every 11-year and curvature and gradient drifts in the large-scale magnetic field of the heliosphere (Jokipii *et al.* 1977; Jokipii & Thomas 1981; Smith 1990).

Recently, features of the interplanetary medium have been explained on the basis of heliospheric neutral current sheet, which separates the whole heliosphere into two regions of opposite polarity of magnetic field. In each hemisphere the field is well approximated by a Parker Archimedian spiral with the sense of the field being outward in one hemisphere and inward in the other. The field direction in each hemisphere altered in each 11-year sunspot cycle. At the solar minimum, the current sheet is nearly

equatorial with the northern hemisphere solar magnetic field being in one direction and the southern magnetic field having the opposite sign. The solar magnetic field structure near the sunspot maxima is complex, where it corresponds roughly to increasing the inclination of the current sheet. The inclinations of the heliosphere neutral current sheet along the equatorial plane of heliosphere are often named as tilt angles. The waviness of neutral current sheet, i.e., tilt angle has been used as solar/interplanetary index by various investigators to explain the long-term modulation of cosmic rays (Webber & Lockwood 1988; Swinson & Yasue 1992; Ahluwalia 1992). The tilt angle is computed by averaging the maximum latitude through the neutral line in the north and south hemispheres in each Carrington rotation. The heliospheric neutral current sheet and its waviness provide us some basic physical mechanism to explain the long-term modulation of galactic cosmic rays.

Many researchers have studied that correlation between CRI and TA is better during $qA < 0$ than $qA > 0$ (Belov 2000; Iskra & Wybraniec 2001; Usoskin *et al.* 2003). In this paper we have made an attempt to correlate CRI with TA and SSN to explain the momentary behaviour of cross correlation function with respect to time (by running cross correlation method) during the whole investigation period. Moreover, the time-lag analysis between CRI and solar activity (SA) parameters has also been performed by the 'minimizing correlation coefficient method' and these results have been verified by the different shapes of hysteresis curves obtained by the cross plot of CRI and SA parameters.

2. Data and method of analysis

In the present work, SSN and TA (waviness of heliospheric neutral current sheet) have been considered as SA parameters to study the long-term modulation of CRI for the period of 1976 to 2005. To study the average behaviour of cosmic ray intensity, monthly mean values of neutron monitor stations of different cut-off rigidity (Oulu, Kiel and Huancayo) have been used, whereas the values of TA were obtained from the Wilcox Solar Observatory (WSO, classical model).

The correlation coefficient between cosmic ray intensity and different solar activity parameters with time lag has also been calculated for the said period using the method of 'minimizing correlation coefficient method'. Here we have selected both the series CRI and TA for the same period with zero time lag and then shifted one series by a step of one month and calculated the cross correlation coefficient between both the series. Similarly, the other series has also been shifted by one month and the new value of cross correlation coefficient is calculated. As such, the time (number of shifted months) is obtained, when the anti-correlation coefficient is maximum. This is the time lag between both the series CRI and TA. The probable error for each value of correlation coefficient has been calculated by the formula: $P.E. = 0.6745(1 - r^2)/\sqrt{N}$, where r is correlation coefficient and N is the number of samples.

In the present paper the 'Running cross correlation method' has been used to study the relationship between CRI and solar activity indices (Usoskin *et al.* 1998; Mishra & Tiwari 2003). In the said method we use a time window of width T centered at time t : $[t - T/2, t + T/2]$. The cross correlation coefficient $c(t)$ is calculated for data within this window. Then the window is shifted in time by a small time step $\Delta t < T$ and the new value of the cross correlation coefficient is calculated. Here we have used the time shifting of one month to calculate the correlation coefficient for

each month between CRI and SSN and for CRI and TA for the period 1976 to 2005. The time window of 50 months has been taken. This value was chosen to match two contradictory requirements; (i) uncertainty of the calculated $c(t)$ is smaller for large T and (ii) T should be small in order to reveal the fine temporal structure of the cross correlation function.

Moreover, the hysteresis curve between CRI and SSN, and CRI and TA has been sketched by taking 30 months moving average of both the data series.

3. Results and discussion

The relationships of sunspot numbers and TA to cosmic ray intensity have been studied earlier (Cliver *et al.* 1996; Cliver and Ling 2001). The inverse correlation between TA and cosmic ray intensity along with 22-year patterns is observed in the evolution of TA. Here an attempt has been made to extend the study to the recent period in order to establish the relationship of sunspot numbers and TA to cosmic ray intensity considering low (Oulu, $R_c \sim 1$ GV), middle (Kiel, $R_c \sim 3$ GV) and high (Huancayo, $R_c \sim 13$ GV) cut-off rigidity neutron monitor stations for the period 1976 to 2005 (solar cycles 21, 22 and 23).

To see the associative behaviour of different cut-off rigidity stations with TA, we have used the % of monthly mean value of CRI for Oulu ($R_c \sim 1$ GV) Kiel ($R_c \sim 3$ GV) and Huancayo ($R_c \sim 13$ GV) from 1976 to 2005. Figure 1 shows overall inverse correlation between TA and % CRI (100% normalized at May 1965) of all the three stations during the whole period of investigation. Looking at the similar behaviour of low to high cut-off rigidity stations, we have chosen the monthly mean value of Kiel ($R_c \sim 3$ GV) a middle cut-off rigidity neutron monitor station. The variation of CRI (Kiel) and TA

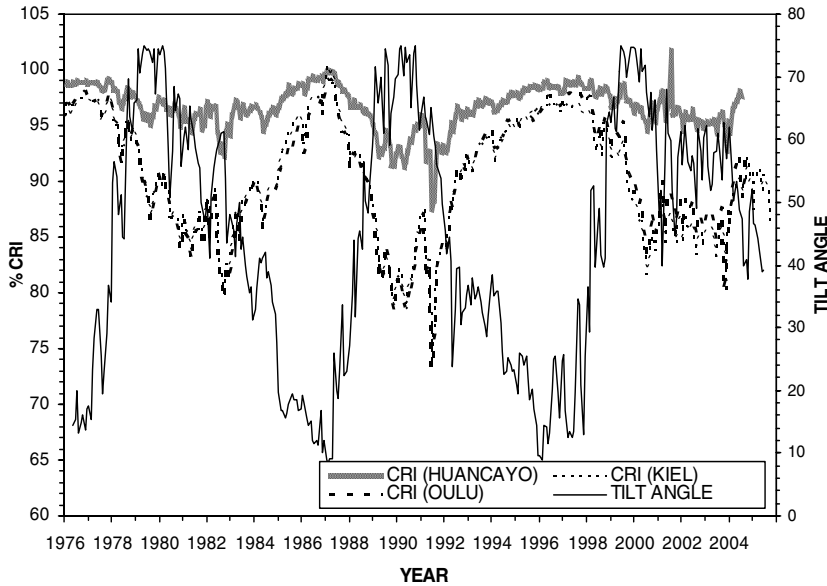


Figure 1. Shows the long-term variation of cosmic ray intensity observed by Oulu, Kiel and Huancayo neutron monitors along with tilt angle from 1976 to 2005.

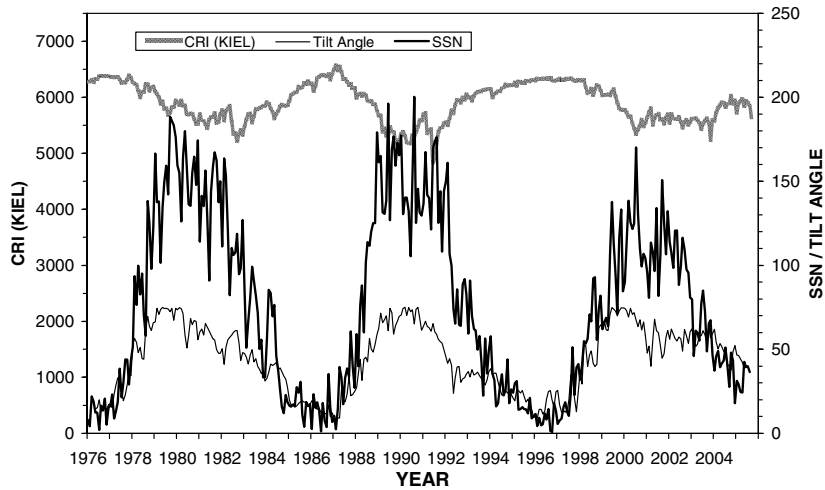


Figure 2. Shows the long-term variation of tilt angle and sunspot numbers with monthly count rates of cosmic ray intensity observed at Kiel neutron monitor for the period 1976 to 2005.

along with sunspot numbers from 1976 to 2005 is shown in Fig. 2. The sunspot number and TA shows a similar pattern and high degree of correlation (positive) with each other whereas cosmic ray intensity is inversely correlated with TA as well as with sunspot numbers with some period time lag during the whole period of investigation. Now we have calculated the cross correlation coefficient between CRI and TA and between CRI and SSN by shifting of both the series one by one by a step of one month. The cross correlation coefficient factor between CRI and TA and between CRI and SSN with different time lags and statistical error bars for solar cycles 21, 22 and 23 is shown in Figs. 3 and 4, respectively. It is observed that during odd cycles 21 and 23 the time lag between CRI and TA is ~ 17 and 11 months at the time of maximum anti-correlation coefficient ($r \sim -0.8$) whereas for even cycle 22 the time lag has been found to be ~ 2 months at the time of maximum anti-correlation coefficient ($r \sim -0.9$). It is ~ 12 and 14 months for odd solar cycles 21 and 23 and ~ 4 months for even solar cycle 22 in the case of CRI and SSN (Table 1). Now we have calculated the running cross correlation between CRI and TA and also for CRI and SSN. From Fig. 5, it is observed that running cross correlation function $c(t)$ is positive during the maxima of odd cycles 21 and 23 for both the cases, i.e., for CRI–SSN and CRI–TA. However, the value of cross correlation coefficient is almost similar in the case of CRI–TA relationship (~ 0.6) for both the cycles 21 and 23 and it is different in the case of CRI–SSN, which is ~ 0.3 and ~ 0.08 for cycles 21 and 23, respectively. This type of analysis is necessary to explain the momentary behaviour of cross correlation function with respect to time, the value of correlation coefficient is different for the different phases of the same solar cycle and it changes with time. The values obtained by this method if averaged over a cycle, will represent the correlation coefficient for a particular cycle. This shows the 22-year variational pattern of cosmic ray intensity and supports the odd–even hypothesis of the CRI cycles. The differences observed in the relationship between CRI and SSN and CRI and TA is perhaps attributable to the different sunspot activity in solar cycles 21 and 23, which is also clear from Fig. 2.

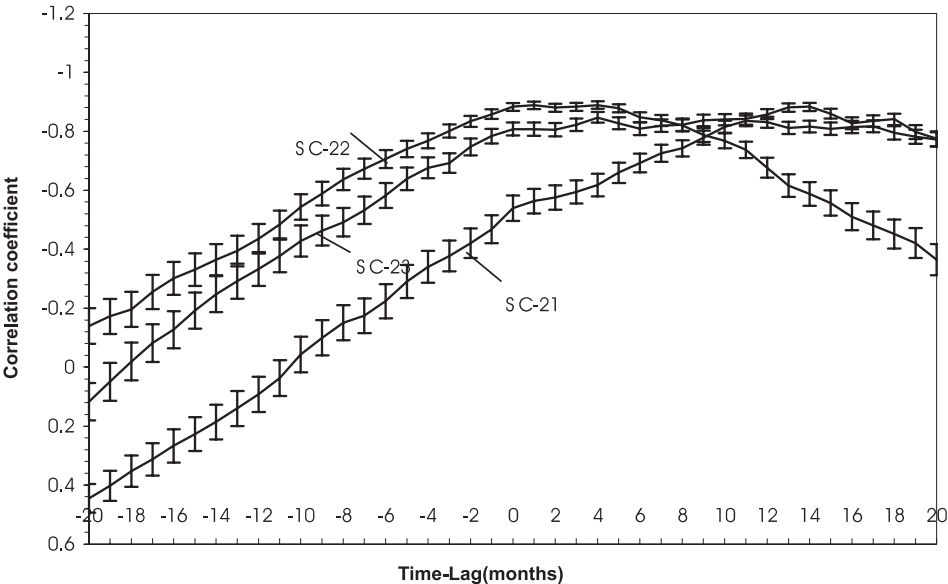


Figure 3. Shows the cross correlation coefficient factor between cosmic ray intensity (Kiel) and tilt angle with different time lags for solar cycles 21, 22 and 23. The statistical error bars are also shown for each value of the correlation coefficient.

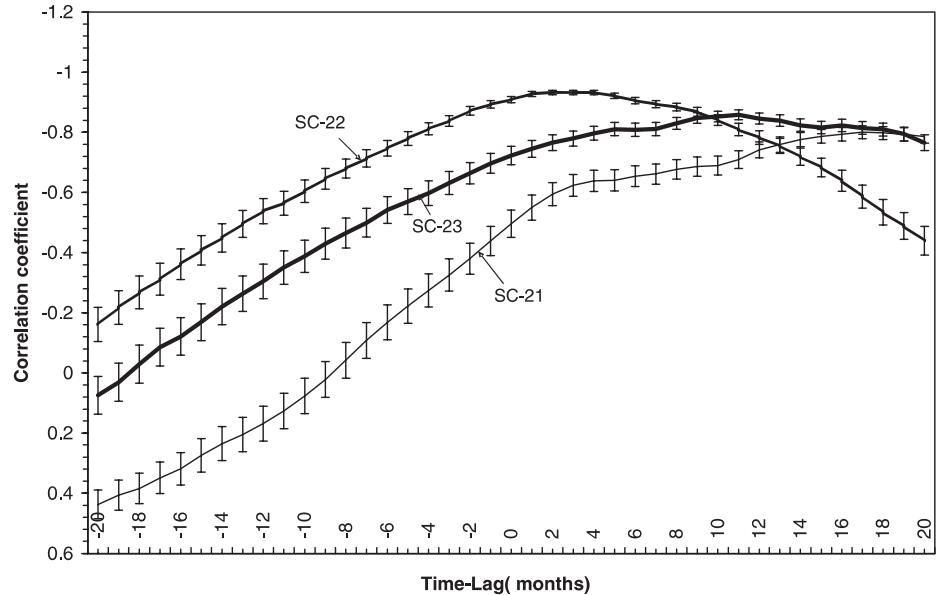
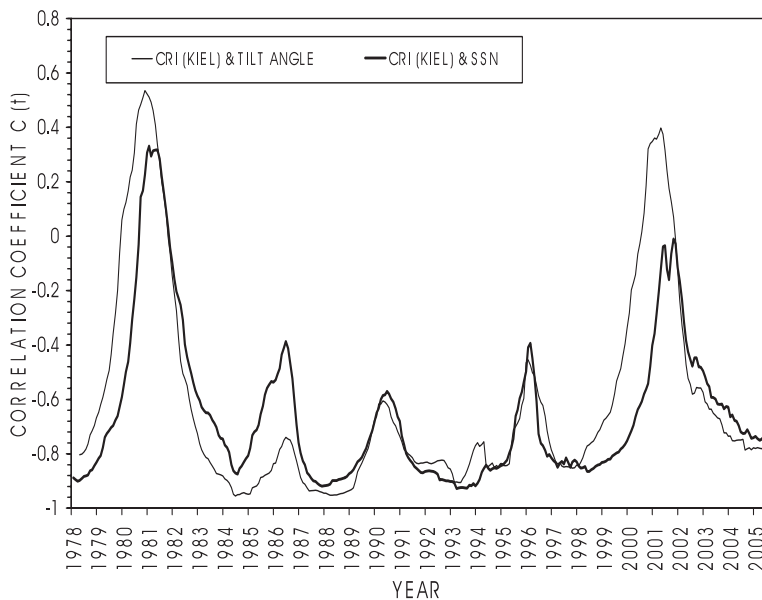


Figure 4. Shows the cross correlation coefficient factor between cosmic ray intensity (Kiel) and sunspot numbers with different time lags for solar cycles 21, 22 and 23. The statistical error bars are also shown for each value of the correlation coefficient.

Table 1. Correlation coefficient between CRI and SSN, and CRI and TA without time lag and with time lag for solar cycles 21, 22 and 23.

Solar cycles	CRI–SSN			CRI–TA		
	Correlation coefficient		Time lag (months)	Correlation coefficient		Time lag (months)
	Without time lag	With time lag		Without time lag	With time lag	
21	−0.539	−0.835	12	−0.495	−0.8002	17
22	−0.882	−0.888	4	−0.908	−0.932	2
23	−0.807	−0.882	14	−0.722	−0.858	11

**Figure 5.** Shows the running cross correlation coefficient $c(t)$ between cosmic ray intensity (Kiel) and sunspot numbers as well as between cosmic ray intensity (Kiel) and tilt angle from 1976 to 2005.

From Fig. 5, it is also evident that there is a 5.5-year periodicity in the observed peaks occurred which is half of the (11-year) solar cycle period. The TA behaviour is similar during the rising phases of the solar cycles 21, 22 and 23 and different during the declining phase of the solar cycle 23 than the solar cycles 21 and 22 (Fig. 2). The similarities in the TA evolution during the rising of cycles 21 and 22 have also been reported (Suess *et al.* 1993; Cliver 1993; Cliver & Ling 2001).

To support the time lag findings, we have further plotted the hysteresis curves between CRI and TA as well as between CRI and SSN, which are shown in Figs. 6 (a, b and c) and 7 (a, b and c), respectively, for the solar cycles 21, 22 and 23. It has been observed that the hysteresis loops for CRI–TA and CRI–SSN are wider for odd cycles and narrow for even cycles, which supports the even–odd asymmetry of the cycles.

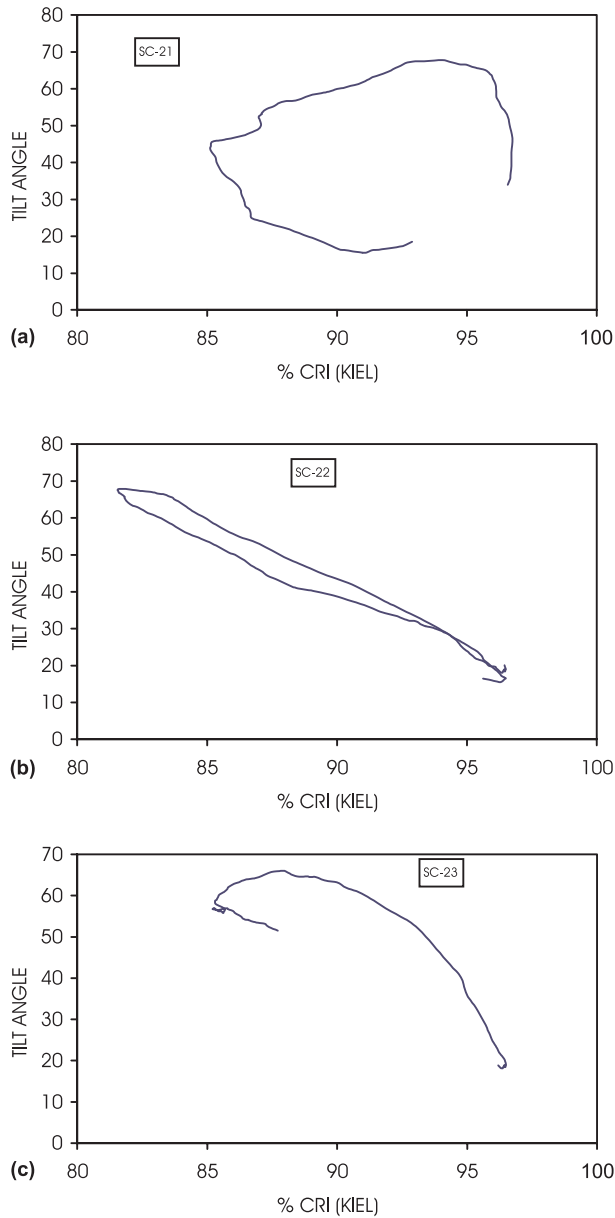


Figure 6 (a, b and c). Shows the hysteresis curves between CRI and tilt angle for the solar cycles 21, 22 and 23, respectively. The different shapes of curves for odd and even cycles are clearly apparent from the figure.

In the drift formulation of cosmic-ray modulation (Jokipii *et al.* 1977; Potgieter 1998), positively charged cosmic rays preferentially enter the heliosphere from the direction of the solar poles during $qA > 0$ cycles (corresponding to times when the polarity of the solar magnetic field is outward in the northern hemisphere) such as ~ 1970 – 1980 and ~ 1990 – 2000 . During $qA < 0$ periods such as ~ 1980 – 1990 when

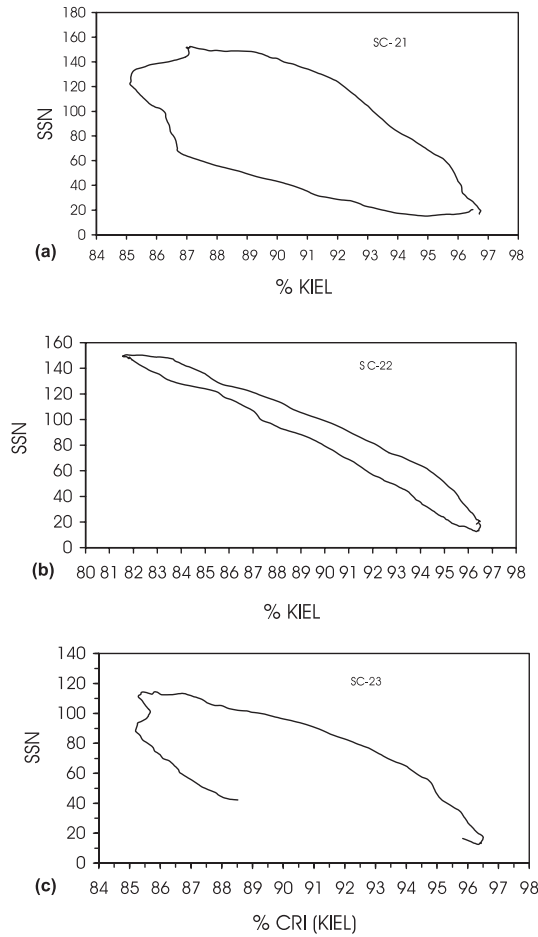


Figure 7 (a, b and c). Shows the hysteresis curves between CRI and SSN for the solar cycles 21, 22 and 23, respectively. The different shapes of curves for odd and even cycles are clearly apparent from the figure.

the solar field polarity is reversed, cosmic rays (positively charged) approach the Sun from along the HCS. During $qA > 0$ times, it might be expected that incoming cosmic rays will be less affected by drift effects associated with an increase in the TA at the beginning of a solar cycle (odd-numbered) or by diffusion associated with enhanced coronal mass ejection (CME) activity. CMEs, which are thought to be a key element in diffusion/convection-based pictures of modulation (Burlaga *et al.* 1984), are characteristically confined to the Sun's equatorial regions early in the solar cycle and appear at higher latitudes during the course of the cycle as the streamer belt at the base of the HCS moves poleward. At the beginning of even-numbered cycles ($qA < 0$), when cosmic rays approach the Sun along the HCS, they will be more readily affected by changes in the TA and low-latitude CMEs. Thus, the difference in the responsiveness to solar activity changes at the onset of even- and odd-numbered solar cycles is consistent with a drift effect (Smith 1990).

While the TA increase was remarkably similar during the rise phase of the last three cycles (Fig. 2), there is evidence that HCS evolution may differ on the decline of even and odd-numbered solar cycles. Specifically, the TA appears to collapse to low angles more rapidly during the decline of even-numbered cycles such as 22 (peak in ~ 1990). We conclude that the differences observed in the relationship between CRI and SSN, and CRI and TA may be due to the low activity of the solar cycle 23.

The understanding of the solar modulation of galactic cosmic rays is still based on the standard model of diffusion, convection and adiabatic deceleration effect, where the interplanetary magnetic field lines including drift processes determine the path of individual particles through the heliosphere. This leads to characteristic differences between adjacent solar cycles due to the different polarity of the solar and large-scale interplanetary magnetic fields. The polarity of the solar magnetic field reverses the sign about every 11-year near the time of maximum solar activity. Thus successive activity maxima are characterized by different solar field polarity. However, for a better understanding of odd-even cycle's differences, the influences of curvature of interplanetary magnetic field on the transport of cosmic ray should also be considered.

Acknowledgements

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