

Editorial

The Physical Research Laboratory (PRL), Ahmedabad celebrated its Diamond Jubilee during 2006–07. An international conference on **Challenges for Solar Cycle-24** was organized during 22–25 January 2007 at PRL, as a part of this celebration. The conference was attended by more than 140 participants including 40 delegates from outside India. The conference was conceived with the idea to address key science issues such as: What would the behaviour and magnitude of the next solar cycle be? What are the major questions to be addressed by the international programs, viz., IHY, Space-Weather, CAWSES and Sun–Earth connection?

An International Advisory Committee (IAC) and Scientific Organizing Committee (SOC) consisting of distinguished scientists from all over the globe were formed to plan the conference. The IAC and SOC decided on three thematic topics, viz., **Magnetic fields, Energetics and Instrumentation**. For each thematic topic, several sub-topics were chosen by the SOC. The topic ‘Magnetic fields’ covered many important subjects such as evolution of magnetic fields, dynamo, helioseismology, helicity, magneto-convection, and coronal magnetic fields. The second topic ‘Energetics’ dealt with sub-topics such as solar irradiance, coronal heating, flares, coronal mass ejections, particle acceleration, IP shocks, solar energetic particles (SEPs), space-weather and Sun–Earth connection, and IHY and CAWSES programs. The ‘Instrumentation’ theme covered ground-based high-resolution observations, active and adaptive optics, next generation instrumentation, current space missions, and future space instruments.

The conference was inaugurated on 22 January 2007 by Prof. U. R. Rao, Chairman, Council of Management, PRL. The inaugural program included a keynote address on *Outstanding Problems in Solar Physics* by Prof. Markus J. Aschwanden of Solar & Astrophysics Laboratory, Lockheed Martin Advanced Technology Center, USA, who gave a broad perspective of the current status of solar physics and its challenges.

The scientific program included plenary, invited and contributory talks. The plenary talks comprised of solicited reviews each of duration 40 min followed by in-depth discussions of about 10 min. A total of six plenary talks were organized: three in ‘magnetic fields’ and three in ‘energetics’. A total of 26 invited talks each of duration 25 min including in-depth discussions were organised covering all three focal topics and addressing the international scenario plus very recent work on a sub-topic of the focal theme topic. A total of 26 oral contributory talks each of duration 15 min were organized in order to provide an opportunity for young researchers as well as for those interested in short presentations and these talks generated very interesting and interactive discussions.

In view of the overwhelming response from various Indian institutes and universities to participate in the conference, we encouraged many participants to make poster presentations. In response, 59 posters from India and outside were displayed which resulted in serious discussions and interactions.

The conference was well received in terms of active participation by a large number of solar researchers including many foreign delegates. The keynote address by Prof. Markus Aschwanden, the plenary talks by Prof. Saku Tsuneta on *First results from Hinode Mission* and by Prof. J. M. Davilla on *First Results from STEREO Mission* as also the several interesting presentations on the prediction of “low versus high amplitude of the next solar cycle 24” and their follow-up discussions were interactive among the participants as well as the media.

The Physical Research Laboratory was founded in 1947 by Dr. Vikram A. Sarabhai, initially, at his residence with research on cosmic rays. The laboratory was later, formally established on November 11, 1947, in the M. G. Science College, Ahmedabad, with support from the Karmkshetra Educational Foundation and the Ahmedabad Education Society. The initial focus was research on cosmic rays and the properties of the upper atmosphere. Research in theoretical physics and radio physics were added later with grants from the Atomic Energy Commission. The history of solar research at PRL began with the founding of the Udaipur Solar Observatory in the middle of Fateh Sagar lake in Udaipur in 1975 and the subsequent taking over by the Department of Space in 1981.

A special evening session was organized in the memory of Prof. Arvind Bhatnagar, founder of Udaipur Solar Observatory, Udaipur, which was chaired by his colleague and friend Prof. J. O. Stenflo. A popular evening lecture on ‘The Inconstant Sun’ was delivered by his colleague Prof. S. M. Chitre, which was attended by the conference participants, students, faculty of PRL and invited guests from Ahmedabad.

A special session on INDO-US collaboration in solar physics was organized at USO, Udaipur on 26 January 2007 and new collaborative efforts for ground as well as space programs were initiated.

We are thankful to the **Indian Academy of Sciences, Bangalore** for publishing this special issue of *Journal of Astrophysics and Astronomy* comprising refereed and accepted papers presented in the conference. Our thanks are due to all the referees for reviewing the papers. We thank members of the International Advisory and Scientific Organizing Committees for suggesting magnificent focal topics and contemporary subjects which helped in framing the scientific program.

The conference was made possible by the generous financial support by Prof. J. N. Goswami, Director, PRL, Ahmedabad. Substantial financial assistance was received from Indo US Science & Technology Forum, New Delhi to organise the special session on ‘INDO-US collaboration in solar physics’ at USO, Udaipur. We also received outstanding support and co-operation from all the members of PRL. We sincerely appreciate the help received from members of the LOC in organizing this conference.

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Guest Editors

Keynote Address: Outstanding Problems in Solar Physics

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Abstract. Celebrating the diamond jubilee of the Physics Research Laboratory (PRL) in Ahmedabad, India, we look back over the last six decades in solar physics and contemplate on the ten outstanding problems (or research foci) in solar physics:

1. The solar neutrino problem
2. Structure of the solar interior (helioseismology)
3. The solar magnetic field (dynamo, solar cycle, corona)
4. Hydrodynamics of coronal loops
5. MHD oscillations and waves (coronal seismology)
6. The coronal heating problem
7. Self-organized criticality (from nanoflares to giant flares)
8. Magnetic reconnection processes
9. Particle acceleration processes
10. Coronal mass ejections and coronal dimming

The first two problems have been largely solved recently, while the other eight selected problems are still pending a final solution, and thus remain persistent *Challenges for Solar Cycle 24*, the theme of this jubilee conference.

Key words. Solar neutrinos—solar interior—helioseismology—solar magnetic field—solar: corona, cycle, flares—magnetohydrodynamics (MHD)—MHD: oscillations, waves—coronal: heating, dimming—self-organized criticality—particle acceleration—coronal mass ejections (CME).

1. The solar neutrino problem

The ultimate source of energy in our Sun is the nuclear burning of hydrogen to helium, which occurs in the central region of the Sun at a temperature of 15 million degrees. This was first understood in the 1920s, when Hans Bethe, George Gamow, and Carl von Weizsäcker identified the relevant nuclear chain reactions, i.e., the proton–proton (p–p) and the carbon–nitrogen–oxygen (CNO) chain reactions. Both chain reactions of nuclear fusion of hydrogen to helium produce neutrinos, which have little interaction with matter and escape the Sun. Solar neutrinos have been detected since 1967 by Raymond Davis Jr. in a chlorine tank in the Homestake Gold Mine in South

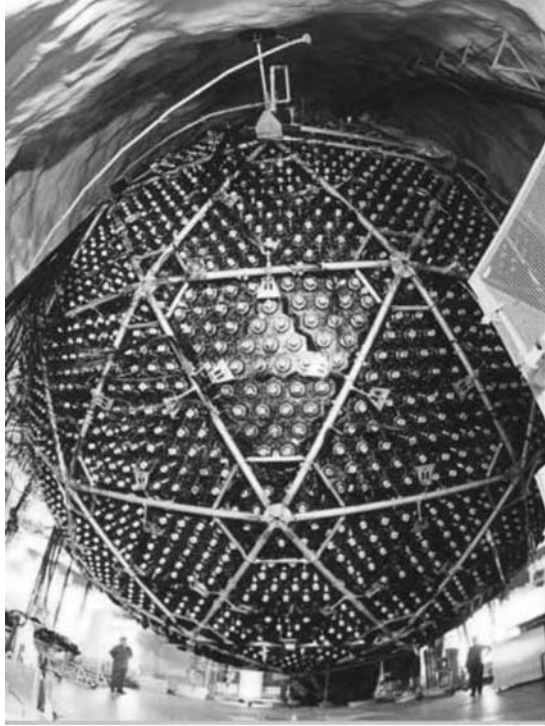


Figure 1. The Sudbury Neutrino Observatory (SNO) in Ontario, Canada, featuring a heavy-water Cherenkov detector that is designed to detect solar neutrinos. It uses 1000 tons of heavy water and 9600 photomultiplier tubes mounted on a geodesic support structure within a 30 meter barrel-shaped cavity, located in the deepest part of the Sudbury mine (Courtesy of SNO).

Dakota, but only about 1/3 of the theoretically predicted neutrino rate was detected over 35 years. This substantial discrepancy raised questions of whether the astrophysical models of the solar interior were incorrect or whether we missed some fundamental part of nuclear physics, which constituted a “power struggle” between astrophysics and particle physics. More powerful particle detectors came online, such as the gallium detector *GALLEX* in Italy and *SAGE* in Russia, sensitive to electronic neutrinos ν_e , as produced by the main nuclear process $4p \mapsto {}^4\text{He} + 2e^+ + 2\nu_e$. However, Pontecorvo and Gribov predicted already in 1969 that low-energy solar neutrinos undergo a *personality disorder* on their travel to earth and oscillate into other flavors of muonic (ν_μ) and tauonic neutrinos (ν_τ), which turned out to be the solution of the *missing neutrino problem*. The first generation of detectors was only sensitive to the highest-energy (electronic) neutrinos, such as the chlorine tank of Raymond Davis Jr, while the *Kamiokande* and *Super-Kamiokande-I* pure-water experiments and the *Sudbury Neutrino Observatory* (SNO, Ontario, Canada; Fig. 1) heavy-water experiments are also somewhat sensitive to the muonic and tauonic neutrinos. So it was the *SNO* that measured for the first time all three lepton flavors (Ahmad *et al.* 2002), which brilliantly confirmed the theory of *neutrino (flavor) oscillations*, and this way also corroborated the astrophysical model and the chemical composition of the solar interior to be

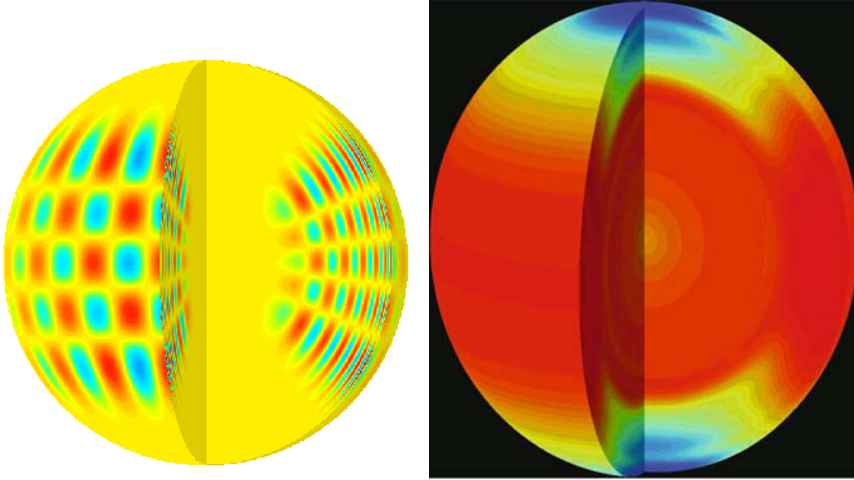


Figure 2. **Left:** A global acoustic p-mode wave is visualized: The radial order is $n = 14$, the angular degree is $l = 20$, the angular order is $m = 16$, and the frequency is $\nu = 2935.88 \pm 0.1 \mu\text{Hz}$, observed with SoHO/MDI (Michelson Doppler Imager). **Right:** The internal rotation rate is shown with a greyscale, measured with SoHO/MDI during May 1996–April 1997. The fastest rotation rate ($P \approx 25$ days) is rendered with a black color, while the slowest rate ($P \approx 35$ days) is given with white color. Note that the rotation rate is rigid in the interior radiative zone, while it is differential in the outer convection zone (Courtesy of SoHO/MDI Team and NASA).

correct (Bellerive 2004; Bahcall 2005). Raymond Davis Jr. garnered the Nobel Prize in Physics 2002 for his pioneering work in neutrino observations (Davis 2003), while the uncrowned John Bahcall, who accomplished most of the fundamental theoretical work around the solar neutrino problem, deceased in 2005.

2. The structure of the solar interior

On one side, the internal structure of the Sun has been inferred from theoretical models: either with hydrostatic equilibrium models, or with time-dependent numerical simulations of the Sun's lifetime evolution. On the other side, experimental verification has been achieved with high accuracy by means of helioseismology, which was discovered in the decade of 1960–1970, from global oscillations on the solar surface in visible light. Velocity oscillations were first measured by R. Leighton, and then interpreted in 1970 as standing sound waves in the solar convection zone by R. Ulrich, C. Wolfe, and J. Leibacher. These acoustic oscillations, also called *p-modes* (pressure-driven waves), are detectable from fundamental up to harmonic numbers of ≈ 1000 , and are most conspicuous in dispersion diagrams, $\omega(k)$, where each harmonic shows up as a separate ridge, when the oscillation frequency (ω) is plotted as a function of the wavelength λ . An example of a p-mode standing wave is shown in Fig. 2 (left). Since the density and temperature monotonically increase with depth inside the Sun, the sound speed varies as a function of the radial distance from Sun center. P-mode waves excited at the solar surface propagate downward and are refracted towards the surface, where the low harmonics penetrate very deep, while high harmonics are confined to the outermost

layers of the solar interior. By measuring the frequencies at each harmonic, the sound speed can be inverted as a function of the depth, and this way the density and temperature profile of the solar interior can be inferred and unknown parameters of theoretical standard models can be constrained, such as the abundance of helium and heavier elements. The detailed elemental abundance of helium and heavier elements is still a key problem of cosmological significance (e.g., Bahcall *et al.* 2005, 2006).

Exploiting the Doppler effect, frequency shifts of the p-mode oscillations can be used to measure the internal velocity rate as a function of depth and latitude, as shown in Fig. 2 (right). A layer of rapid change in the internal rotation rate was discovered this way at the bottom of the convection zone, the so-called *tachocline* (at 0.693 ± 0.002 solar radius, with a thickness of 0.039 ± 0.013 solar radius). The physical structure of the tachocline is still considered as a key to understand the solar dynamo and the generation of magnetic fields (e.g., Gilman 2000).

Besides the p-mode waves, *gravity waves* (*g-modes*) are suspected to exist in the solar core, where buoyancy rather than pressure supplies the restoring force. These gravity waves were predicted to have long periods (hours) and very small velocity amplitudes, but they have not yet been convincingly detected, which represents another open problem about the deepest core of our star (e.g., Rogers & Glatzmaier 2005).

While *global helioseismology* detects p-modes as a pattern of standing waves that encompass the entire solar surface, local deviations of the sound speed can also be detected beneath sunspots and active regions, a diagnostic that is called *local helioseismology* (for reviews see, e.g., *Solar Physics*, Vol. **192**). Near sunspots, p-modes are found to have oscillation periods in the order of 3 minutes, compared to 5 minutes in active region plages and quiet-Sun regions. So, local helioseismology allows us to probe flow patterns beneath sunspots, which yields us clues on magneto-convection and large-scale flows (e.g., meridional flows) associated with the solar dynamo in general. Disturbances of the sound speed can even be used to probe the far-side of the Sun, which enables us to make frontside predictions of active regions that emerge on the backside of the Sun.

3. The solar magnetic field

There are many problems related to the physical understanding of the solar magnetic field, subject of numerous and intense studies. The most fundamental aspect is the workings of the solar dynamo in the solar interior, which generates the magnetic field we observe on the solar surface and in the corona. The basic idea of the solar dynamo has been developed in the work of Parker, Babcock, Leighton, Deubner, and Ulrich in the years after 1955. The solar dynamo consists of a poloidal field at the beginning of the solar cycle that becomes wound up (omega effect) and twisted (alpha effect) progressively due to the differential solar rotation, and finally evolves into a highly stressed toroidal field, which then breaks up by many local reconnection events (associated with flares and coronal mass ejections), leading to the decay of the solar cycle. Although we understand the basic mechanism, we still do not understand a lot of physical aspects of the solar dynamo (e.g., Charbonneau 2005). For instance, analytical models of the dynamo produce too much twisting and thus predict shorter cycles than the observed 11 years. Meridional flows have been detected with helioseismology, slow surface flows of order $\approx 20 \text{ m s}^{-1}$, which seem to play a critical role for the solar dynamo. Moreover, opposite magnetic helicity is observed in weak and strong field

regions, another puzzle that needs to be explained by a successful dynamo model. We do not expect that a fully analytical model will reproduce the solar dynamo, since also chaotic behavior and intermittency are expected to play a role, which make a fully deterministic model problematic.

The solar photosphere is the place where the solar magnetic field is most directly observed (by means of the Zeeman effect), but this is also the magnetically most complex region due to the subphotospheric magneto-convection, which creates the granular, mesogranular, and supergranular structures. The solar dynamo is thought to produce magnetic fibrils in the tachocline and convection zone, which buoyantly rise to the surface, become mixed by turbulent magneto-convection, and become better understood only now with numerical MHD simulations (e.g., Carlsson & Stein 1997) and impressive high-resolution observations (e.g., DePontieu *et al.* 2007).

Besides the generation of the magnetic field in the solar interior, we have also fundamental problems to model the magnetic field in the solar atmosphere and corona. Potential (source-surface) models yield a reasonable first approximation of the coronal field in most cases, but they cannot reproduce twisted field lines (filaments, magnetic flux ropes) or current sheets (important for the formation of flares and coronal mass ejections), because potential fields ignore the currents by definition. Force-free models stand on a physically more sound basis, but the simplest version (of linear force-free fields) renders oversimplified models, while more general versions (nonlinear force-free fields) require large computational efforts and are insufficiently constrained by photospheric magnetograms or magnetograph data (Schrijver *et al.* 2006). There is also the fundamental problem that any extrapolation of photospheric magnetic fields bears significant uncertainties because the lowest 500 km of the chromosphere is not force-free (Metcalf *et al.* 1995). Furthermore, most magnetic field models ignore the density and temperature of the plasma, except models computed with full MHD codes. An observational constraint that is used more and more is the tracing of coronal EUV loops, because they are assumed to trace out exactly the coronal field due to the low plasma- β parameter of the corona (Fig. 3). However, the 3D reconstruction of EUV loops is still in its infancy, but automated loop recognition algorithms and stereoscopic 3D reconstructions with the new STEREO mission are rapidly developing (e.g., Wiegmann & Inhester 2006).

4. Hydrodynamics of coronal loops

Standard models of the solar corona assume gravitational stratification and spherical symmetry, which provides a 1D model of the average density for the dominant temperature as a function of altitude. However, the solar corona is highly inhomogeneous, structured by loops (closed field lines) and plumes (open field lines in coronal holes), and more complex transient structures, which are not necessarily in hydrostatic equilibrium. The most widely used hydrostatic model of coronal loops assumes an equilibrium between a spatially uniform heating rate and the radiative and conductive loss rates, the so-called Rosner–Tucker–Vaiana (RTV) model (Rosner *et al.* 1978). Although this model predicts a scaling law between the loop apex temperature, loop length, and pressure, that approximately corresponds to the observations of hotter soft X-ray emitting loops (Rosner *et al.* 1978), the new EUV observations of coronal loops revealed density deviations up to two orders of magnitude, super-hydrostatic density scale heights, near-isothermality along the loops, and obvious dynamical behaviors,

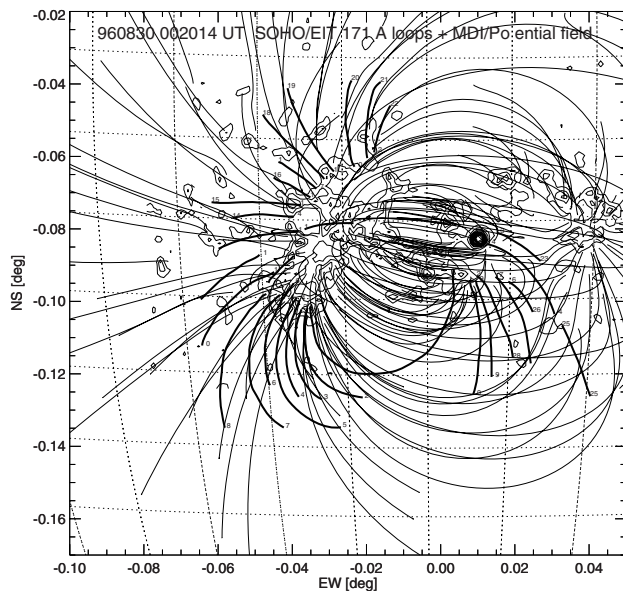


Figure 3. Magnetic field lines of an active region are shown, based on a theoretical potential field model extrapolated from magnetograms observed with SoHO/MDI on August 30, 1996 20:48 UT (thin lines), and compared with tracings of EUV loops observed simultaneously with SoHO/EIT at 171 Å (thick lines). Note the misalignment of cospatial loops and field lines, which indicates the inappropriateness of a potential field model (from Aschwanden *et al.* 1999).

properties that are all not consistent with the RTV model. We can understand that hotter soft X-ray loops agree better with the RTV model than the cooler EUV loops, because the pressure scale height in hotter loops is generally longer than the loop height, and thus the assumption of a constant pressure made in the RTV model is better fulfilled. The cooler EUV loops have a shorter density or pressure scale height, say $\lambda \approx 47,000$ km for 1 MK loops, and thus the gravitational stratification has to be included in hydrostatic models, which has been considered in the model of Serio *et al.* (1981).

However, the least known parameter is the spatial heating function. Serio *et al.* (1981) parameterized the heating function with an exponential heating scale height, which predicted higher loop densities and flatter temperature profiles than the RTV model, as it was actually observed in EUV loops with the TRACE spacecraft (e.g., Lenz *et al.* 1999; Aschwanden *et al.* 2000). Some active region loops were observed to have density scale heights up to four times the hydrostatic thermal scale height (Fig. 4), so they were clearly not in hydrostatic equilibrium. Although difficult to measure, plasma flows were readily detected in a number of loops, which require dynamic models, such as siphon flow models. However, what made hydrostatic models even more questionable was the observed temperature evolution, i.e., EUV loops were clearly observed most of the time to be cooling through the EUV passbands, while the hotter counterpart appeared in soft X-rays typically an hour earlier (Winebarger & Warren 2005). Generally, the observed cooling time was found to be significantly longer than the theoretically expected radiative cooling time, which required hydrodynamic models

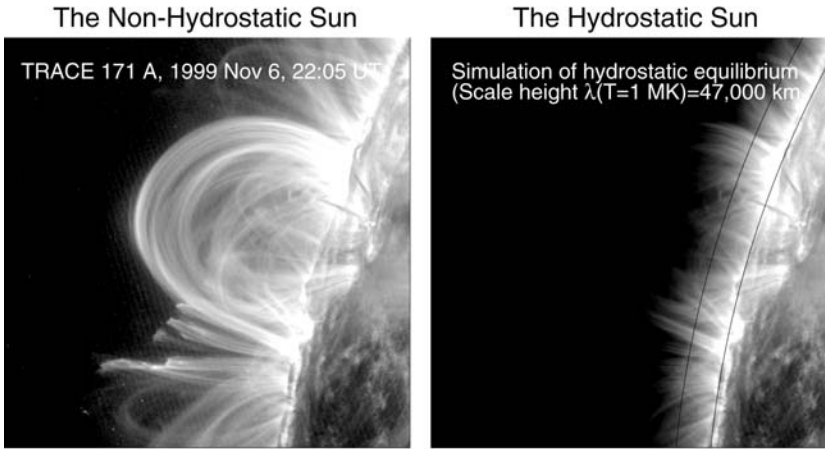


Figure 4. An active region at the limb, observed with TRACE on November 6, 1999, 22 UT, is shown on the left. These loops have a density scale height of at least four times the hydrostatic scale height (as simulated in the right panel), and thus are in a state far away from hydrostatic equilibrium (Aschwanden *et al.* 2001).

with temporally intermittent heating functions (Warren & Winebarger 2006). So, we find ourselves in abandoning hydrostatic models and delving into more complex hydrodynamic models to reconcile the observed density and temperature structure of coronal loops, not to talk about intricate data analysis problems how to separate an individual coronal loop from the complex background that contains myriads of competing other loop structures.

5. MHD oscillations and waves

Much like the discovery of helioseismology four decades ago, it was recently discovered that the solar corona contains an impressively large ensemble of plasma structures that are capable of producing sound waves and harmonic oscillations. Coronal seismology was mostly enabled by the high spatial resolution, image contrast, and time cadences of the SoHO and TRACE spacecraft in EUV images. Oscillating loops (Fig. 5), prominences, or sunspots, as well as propagating waves have been intensively studied since about 1999. While the theory of MHD oscillations was ready since about the 1980s, observational verification occurred only about two decades later. Essentially all predicted MHD wave modes have been observed and verified today, which includes MHD fast mode standing waves (kink modes and sausage modes) and MHD slow mode (acoustic) standing waves, as well as propagating MHD waves. Most of the loop oscillations occur after a filament eruption, flare, and/or coronal mass ejection (CME). It occurs that mass transport during an eruptive phenomenon causes a pressure imbalance in the surrounding coronal plasma, which displaces loops and excites this way oscillations until a new pressure equilibrium is reached. A surprising observation was also that most of these loop oscillations are strongly damped, but there is much theoretical argument about the most relevant damping mechanism. For fast MHD oscillations the two most likely damping mechanisms are phase mixing and resonant

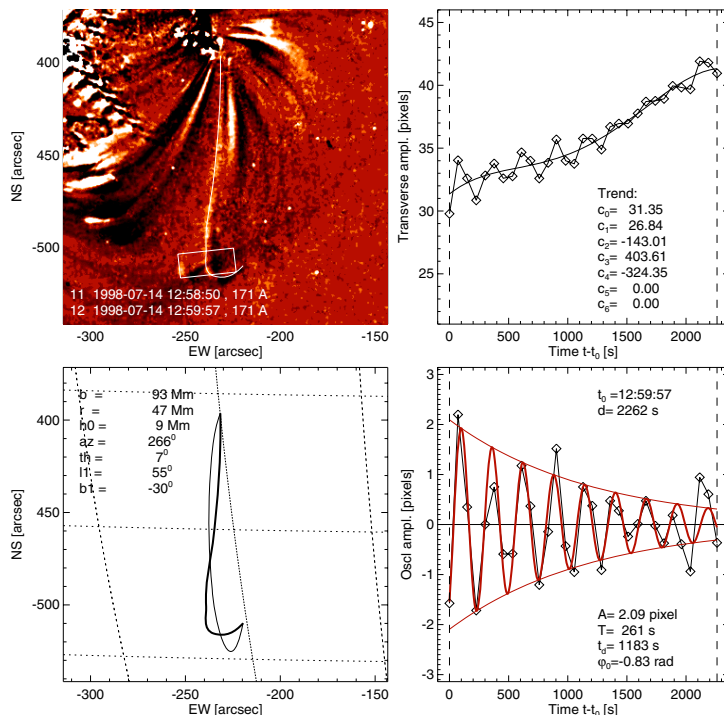


Figure 5. A loop oscillation is observed with TRACE on July 14, 1998, 13 UT: The difference image shows loop displacements with respect to an image taken 1 minute earlier (top left), the coordinates of the oscillating loop show deviations from a semi-circular geometry (bottom left), the loop amplitude measured in perpendicular direction shows also a superimposed trend of centroid motion (top right), and the oscillatory component exhibits strong damping (bottom right) [from Aschwanden *et al.* 2002].

damping, while the slow-mode MHD oscillations seem to be damped by radiative cooling.

Although the young new field of coronal seismology confirmed the theoretically expected MHD modes, there are still a number of significant problems to be worked on, such as the nature of the excitation mechanism, the detailed 3D modeling of asymmetrically excited modes, and the physical nature of the damping mechanism. Alfvénic waves seem also to be excited in the outflows of reconnection regions as reported in some flares. In principle, observations of fast MHD kink-mode oscillations allow us to measure the magnetic field in the solar corona directly, but an accurate determination requires also measurements of the density contrast inside and outside of the loops. One MHD oscillation mode that has not yet been successfully detected is the torsional mode. Detailed 3D modeling of loop oscillations have also to consider the effects of loop curvature, gravitational stratification, the loop cross-section geometry, the polarization direction, coupling of modes, higher harmonics, and temperature effects from radiative and conductive cooling. For recent reviews see, e.g., Aschwanden (2004), Nakariakov & Verwichte (2005).

6. The coronal heating problem

The coronal heating problem has been with us since 6 decades, since Bengt Edlén and Walter Grotrian identified the highly ionized Fe IX and Ca XIV lines in the solar spectrum in 1943, which indicated a coronal temperature of ≈ 1 MK that was about 200 times higher than the underlying chromosphere. Such a hot corona can obviously not be generated by thermal conduction, but requires a special mechanism that turned out to be difficult to detect directly. Today, we have a bifurcation of two mechanisms that seem to be responsible for heating of the solar corona: (1) a *basal heating mechanism* that is most likely related to small-scale reconnection processes in the transition region, which fills coronal loops in active regions and in the quiet Sun, and (2) an *extended corona heating mechanism* that is most likely generated by ion-cyclotron heating of Alfvén waves, which heat the extended corona and solar wind in a typical distance of $\approx 2\text{--}5$ solar radii.

Recent observations show no evidence for local (*in-situ*) heating in the solar corona in active regions and in the quiet Sun, but rather for a basal heating mechanism that operates below the corona in the transition region and upper chromosphere, with subsequent *chromospheric evaporation* as known in flares. New observational evidence for this scenario comes from (1) the temperature evolution of coronal loops, (2) the over-density of hot coronal loops, (3) upflows in coronal loops, (4) the Dopplershift in coronal loops, (5) upward propagating waves, (6) the energy balance in coronal loops, (7) the magnetic complexity in the transition region, (8) the altitude of nanoflares and microflares, (9) the cross-section of elementary loops, as well as from (10) 3D MHD simulations of coronal heating (such as by Gudiksen & Nordlund 2005). The phrase *coronal heating problem* is therefore a paradoxical misnomer and should rather be addressed as *chromospheric heating problem* and *coronal loop filling process*. This paradigm shift substantially reduces the number of relevant theoretical models for coronal heating in active regions and the quiet Sun (Aschwanden 2007).

In contrast, the heating of coronal holes and the solar wind seems to be accomplished by a different mechanism. SoHO/UVCS has revealed surprisingly large temperatures, outflow speeds, and velocity distribution anisotropies for positive ions in coronal holes. The role of Alfvén waves (their generation, propagation, and reflections) becomes more important on larger spatial scales (since Alfvén waves have a long dissipation length of several solar radii). Wave-particle interactions, such as ion-cyclotron resonance, are considered now as the principal mechanism for heating of coronal holes, and ultimately driving the fast solar wind (Hollweg 2006; Cranmer 2002, 2004). The current understanding is that the solar wind is mainly driven by the pressure of hot protons, so the heating in coronal holes goes more into protons than electrons, because it is conveyed by the ion-cyclotron resonance rather than by currents, which is different from the DC heating models generally applied in the lower corona (basal coronal heating).

7. Self-organized criticality in solar flares

Self-organized criticality is universally manifested by powerlaw distributions of system size parameters, measured on spatial, temporal, or energy scales. Powerlaws are the statistical distributions that result from nonlinear processes, while exponential (or Poisson) distributions are the characteristics of random processes with linear scaling.

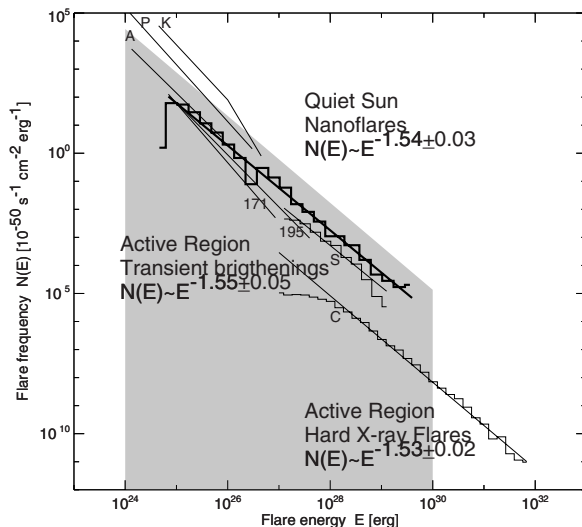


Figure 6. Compilation of frequency distributions of flare energies in the nanoflare range in EUV, active region transient brightenings in soft X-rays and hard X-ray flares. The grey area marks the required energy for coronal heating, which is about a factor of 3 higher than the observed radiated energy.

The principle of self-organized criticality, known from sandpiles, snow avalanches, and earthquakes, has been first applied to solar flares by Lu & Hamilton (1991). Today, the powerlaw found first for large flares, has been extended to microflares (in hard X-rays), active region transient brightenings (in soft X-rays), to nanoflares (in EUV), covering some 8 decades in energy ($\approx 10^{24} - 10^{32}$ erg) (Fig. 6). A crucial question is whether the power law slope at low energies is steeper than the critical value of 2, which would imply that coronal heating is dominantly accomplished by nanoflares and picoflares at the low end of the distribution. At large energies, flares seem to obey a slope of ≈ 1.5 , so the total amount of dissipated energy in the solar corona is dominated by the largest giant flares. At the low end, powerlaw slopes for the smallest detected EUV flares have been reported in a range of ≈ 1.6 up to 2.6, but the steeper slopes seem to be biased by incomplete temperature coverage, as well as by unrealistic geometric models that do not include the proper fractal scaling of flare volumes (Aschwanden & Parnell 2002). Nevertheless, also stellar flares seem to show a range of slopes from ≈ 1.5 to ≈ 2.5 (Audard *et al.* 2000), which are not yet properly understood.

8. Magnetic reconnection processes

Although everybody agrees that magnetic reconnection processes are the principal drivers of all eruptive phenomena on the Sun, such as eruptive filaments and prominences, flares, and coronal mass ejections (CME), the detailed magnetic topologies remain subject of intense investigations and may substantially differ from what we can observe *in situ* in the Earth's magnetosphere. The main observational handicap is the lack of illumination and insufficient spatial resolution to map out the magnetic pre-reconnection configuration, while the post-reconnection configuration is usually

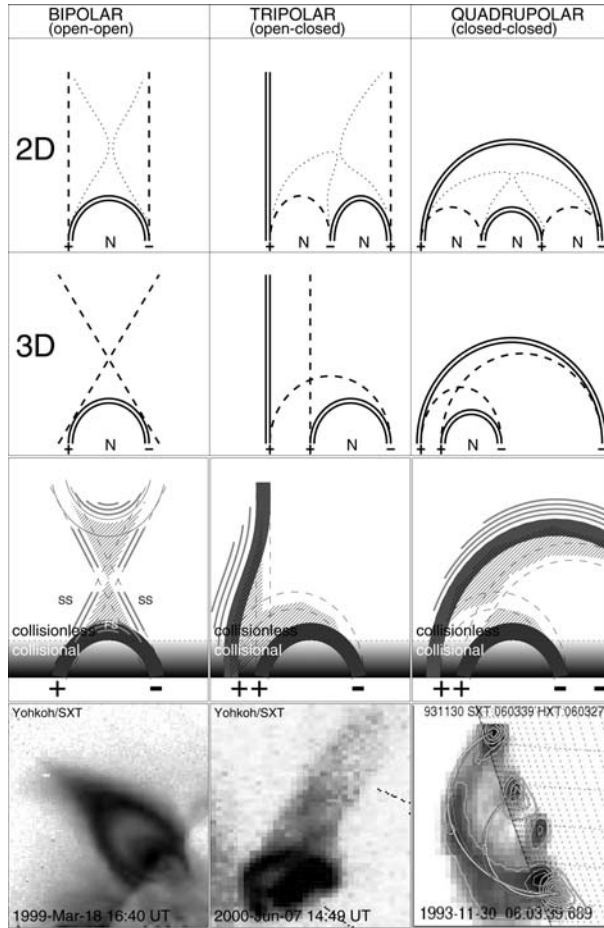


Figure 7. The topology of magnetic reconnection regions is classified into three classes: bipolar or open-open (left column), tripolar or open-closed (middle column), and quadrupolar or closed-closed field line reconnection (right column). The 2D versions are shown in the top row, with the pre-reconnection field lines marked with dashed lines, during reconnection with dotted lines, and post-reconnection field lines with double solid linestyle. The 3D versions are indicated in the second row, where the pre-reconnection field lines are not coplanar, but located behind each other. The third row indicates the acceleration regions (hatched), the relative densities (grey scale), and upward/sideward directed shocks (grey lines). The bottom row shows flare observations from Yohkoh/SXT that correspond to the three different reconnection topologies.

prominently illuminated in soft X-rays and EUV from the chromospherically evaporated hot plasma that fills the post-flare loop system. Therefore, we have to reconstruct the pre-reconnection configuration from topological constraints. The simplest basic configurations can be classified as bipolar, tripolar, and quadrupolar configurations (Fig. 7), which can be approximated by 2D models, or more realistically, with 3D models (Fig. 7, top two rows). Bipolar flares produce also ejecta and Masuda-type above-the-loop-top hard X-ray sources; tripolar flares produce also linear soft X-ray jets; and quadrupolar flares have a stricter confinement of the flare plasma and

might even have no access to open field lines, and thus might lack radio emission from escaping electron beams (i.e., type III bursts). In the meantime we found many more flare by-products that allow us to study quantitatively the process of magnetic reconnection, such as: reconnection inflows detected in EUV, reconnection outflows detected in EUV, footpoint motion and the magnetic field in the flare ribbons that allow us to quantify the evolution of the reconnection rate, high-temperature ridges detected in soft X-rays that mark slow shocks and fast-mode standing shocks predicted by the Petschek reconnection model, and bi-directional electron beams detected in radio and hard X-rays that reveal the symmetry of X-points. More complex scenarios in terms of 3D reconnection have also been inferred from observations, indicating intriguing hints of fan, dome, or spine reconnection.

But there still remain a number of open issues on magnetic reconnection in solar flares:

- the initial configuration and driver (unstable filament or prominence, photospheric flux emergence, converging flows, or shear flows);
- the origin of anomalous resistivity in current sheets and the observed intermittent time structures, which seem to indicate a bursty reconnection mode rather than steady Sweet-Parker or Petschek type reconnection;
- the frequent lack of and low value of inflow velocities observed so far;
- the rare detection of outflows and their dynamic nature, which sometimes is observed to produce oscillatory waves;
- the open-field component at the onset of reconnection that is essential to produce the observed type III bursts;
- the electron number problem between the available coronal flare volume and the hard X-ray producing energetic particles, as well as the high asymmetry in the number of upward and downward accelerated particles. For recent textbooks that deal with magnetic reconnection in solar flares see, e.g., Priest & Forbes (2000); Tajima & Shibata (2002); Aschwanden (2004).

9. Particle acceleration processes

There are three main categories of particle acceleration processes that are thought to play an important role in solar flares and CMEs:

- electric DC-field acceleration,
- stochastic acceleration, and
- shock acceleration (for reviews see, e.g., Miller *et al.* 1997; Aschwanden 2004).

We have realistic scenarios for all three mechanisms and believe that they all can operate in solar flare configurations. The electromagnetic fields near magnetic X-points can accelerate particles in current sheets, wave turbulence provides a suitable environment for stochastic acceleration, and the standing fast-mode shocks in reconnection outflows or the shocks at the fronts and flanks of CMEs are all capable of producing shock acceleration. While there is no shortage of theoretical models, our main problem in understanding which acceleration mechanism is most relevant in solar flares is an observational one. The acceleration mechanism itself is a black box for our instruments, while we can only observe secondary emissions such as hard X-ray bremsstrahlung

once the nonthermal particles slam into the chromosphere, or radio emission once electron beams are formed and propagate into interplanetary space. A key question is also whether solar energetic particles (SEP) observed at 1 AU originate in the coronal flare site or in interplanetary CME shocks. Even detailed energy-dependent timing and spectral studies do not tell us first-hand information about the acceleration mechanism, because the secondary signatures represent a convolved signal of at least 5 different processes: acceleration, injection, propagation, trapping, and energy loss. What is needed to make progress is to measure the 3D magnetic topology in the reconnection region and in shocks, the 3D reconstruction of the electromagnetic fields, the 3D reconstruction of the particle trajectories, and physical modeling of the collisional plasma along their trajectories and energy loss sites – not a small task!

10. Coronal mass ejections

Coronal mass ejections (CMEs) are the last phenomenon in the chain of magnetic instabilities in the solar corona, which start with a filament eruption, generate a flare, and launch a CME (for reviews see, e.g., Alexander 2006; Gopalswamy *et al.* 2005; Zhang & Low 2005; Bothmer & Daglis 2007). CMEs received attention only in the 1980s, when the white-light coronagraphs of SMM showed that CMEs are as frequent as flares. Today, with the coronagraphs LASCO on SoHO and on STEREO, the golden age of CME research started. The chief problems we study are:

- their origin, which is often the same magnetic instability that is responsible to trigger filament eruptions and flares,
- the evolution and propagation of CMEs in interplanetary space, including their spatial topology (twisted flux ropes or half shell-like bow-shocks),
- the dynamic coronal response to a CME launch (coronal dimming, EIT waves, loop oscillations),
- the hydrodynamic modeling of CMEs (in terms of MHD codes that simulate the propagation of a CME shock and its interaction with the ambient solar wind),
- the capability, timing, geometry, and energetics of particle acceleration in CME shocks, and
- the geoeffectivity of CMEs and related space weather effects, a catalog of research problems that becomes bigger every year.

11. Final remarks

This brief sketch of ten outstanding problems in solar physics illustrates what a vibrant research field has been opened up, thanks to the multi-wavelength observations that became available to us by solar-dedicated space missions over the last six decades, such as: Skylab (1973), SMM (1980–1989), Yohkoh (1992–2001), Compton GRO (1991–2000), SoHO (1995–today), TRACE (1998–today), RHESSI (2002–today), and SOXS (2003–today). Now, this golden era is continuing to thrive with new missions such as STEREO (launched in October 2006), HINODE (launched in August 2006), and SDO (to be launched in 2008). We note also that a fundamental transformation took place in our approach to study the Sun over the last six decades: we started from the observationally dominated *solar astronomy* and evolved now into true *solar physics*, where a deeper physical understanding is accomplished by theoretical and numerical modeling.

Acknowledgements

I warmly thank Rajmal Jain and the Organizing Committee of this Jubilee Conference for the generous hospitality at PRL, and Arnab Choudhary for showing me a fascinating tour through Ahmedabad and the Gandhi Ashram in Gujarat. This work is partially supported by NASA contract NAS5-98033 of the RHESSI mission through University of California, Berkeley (subcontract SA2241-26308PG), NASA contract NAS5-38099 for the TRACE mission, and NASA contract N00173-02-C-2035 for the STEREO mission through NRL.

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Solar Magnetic Fields

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Abstract. Since the structuring and variability of the Sun and other stars are governed by magnetic fields, much of present-day stellar physics centers around the measurement and understanding of the magnetic fields and their interactions. The Sun, being a prototypical star, plays a unique role in astrophysics, since its proximity allows the fundamental processes to be explored in detail. The PRL anniversary gives us an opportunity to look back at past milestones and try to identify the main unsolved issues that will be addressed in the future.

Key words. Sun—magnetic fields—polarimetry—history of science.

1. Looking back at the past century

I wish to dedicate this presentation to my dear friend Arvind Bhatnagar, whom I have known for nearly four decades, since we shared office for half a year in 1968 in Pasadena, California. Arvind later founded the Udaipur Solar Observatory, which has played an important role both for Indian science and as a partner in international projects. While remembering Arvind and celebrating the Diamond Jubilee of six decades of PRL, let me briefly recall some key personalities and events of the past century that are related to solar magnetic fields.

On the time scale of PRL, now more than six decades ago, the Swedish physicist Hannes Alfvén published a brief letter to *Nature* (Alfvén 1942, the year I was born), in which he tried to explain the origin of sunspots. A byproduct of this theory was his introduction of magnetohydrodynamic waves. Almost three decades later, in 1970, Alfvén received the Nobel prize for this, although his sunspot theory may be considered as incorrect. Many of us can take heart from this example: Do not be afraid to publish a theory that may turn out to be wrong, because it might contain some element of lasting value that may lead to a Nobel prize! In Fig. 1, Alfvén’s brief, Nobel-prize winning paper is reproduced, together with a photo that I took of him (in guru position) with his wife in 1968 (the year I met Arvind, and two years before Alfvén’s Nobel prize), in front of their home in La Jolla, California.

Let us look back further. More than eight decades ago, in 1923, Wilhelm Hanle in Göttingen discovered what is now called the Hanle effect (Hanle 1924), which shows how magnetic fields break the symmetry of coherently superposed quantum states (Schrödinger cat states) and cause partial decoherence that increases with the strength of the field. This discovery played a key role in clarifying and understanding the central concept of linear superposition of quantum states in the early days of quantum



Existence of

Electromagnetic-Hydrodynamic Waves

If a conducting liquid is placed in a constant magnetic field, every motion of the liquid gives rise to an E.M.F. which produces electric currents. Owing to the magnetic field, these currents give mechanical forces which change the state of motion of the liquid. Thus a kind of combined electromagnetic-hydrodynamic wave is produced which, so far as I know, has as yet attracted no attention.

The phenomenon may be described by the electrodynamic equations

$$\text{rot } H = \frac{4\pi}{c} i$$

$$\text{rot } E = -\frac{1}{c} \frac{dB}{dt}$$

$$B = \mu H$$

$$i = \sigma \left(E + \frac{v}{c} \times B \right);$$

together with the hydrodynamic equation

$$\frac{dv}{dt} = \frac{1}{\rho} (i \times B) - \text{grad } p,$$

where σ is the electric conductivity, μ the permeability, ρ the mass density of the liquid, i the electric current, v the velocity of the liquid, and p the pressure.

Consider the simple case when $\sigma = \infty$, $\mu = 1$ and the imposed constant magnetic field H_0 is homogeneous and parallel to the z -axis. In order to study a plane wave we assume that all variables depend upon the time t and z only. If the velocity v is par-

allel to the x -axis, the current i is parallel to the y -axis, and produces a variable magnetic field H' in the x -direction. By elementary calculation we obtain

$$\frac{d^2 H'}{dz^2} = H_0^2 \frac{d^2 v}{dz^2},$$

which means a wave in the direction of the z -axis with the velocity

$$V = \frac{H_0}{\sqrt{4\pi\rho}}.$$

Waves of this sort may be of importance in solar physics. As the sun has a general magnetic field, and as solar matter is a good conductor, the conditions for the existence of electromagnetic-hydrodynamic waves are satisfied. If in a region of the sun we have $H_0 = 16$ Gauss and $\rho = 0.005$ gm. cm.⁻³, the velocity of the waves amounts to

$$V \sim 60 \text{ cm. sec.}^{-1}.$$

This is about the velocity with which the sunspot zone moves towards the equator during the sunspot cycle. The above values of H_0 and ρ refer to a distance of about 10^4 cm. below the solar surface where the original cause of the sunspots may be found. Thus it is possible that the sunspots are associated with a magnetic and mechanical disturbance proceeding as an electromagnetic-hydrodynamic wave.

The matter is further discussed in a paper which will appear in *Arkiv för matematik, astronomi och fysik*.

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Figure 1. Hannes Alfvén's Nobel-prize winning *Nature* paper from 1942. At upper left, he sits with his wife outside their home in La Jolla, summer of 1968.

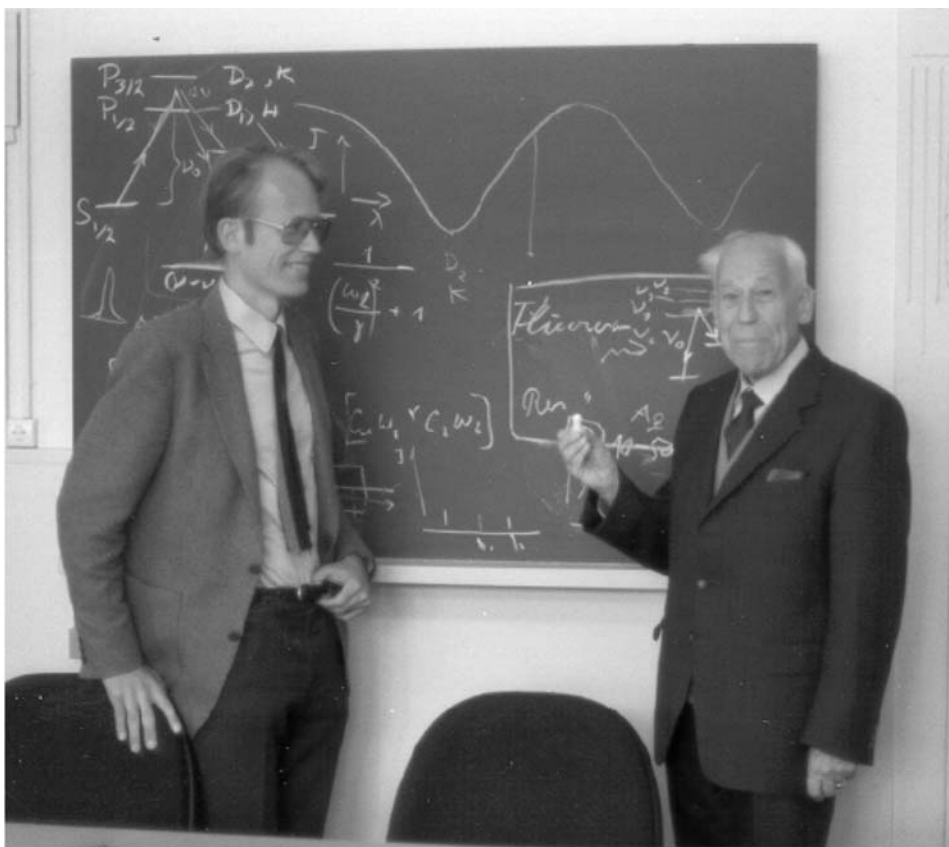


Figure 2. Wilhelm Hanle in my office at ETH Zurich in 1983, on the occasion of the 60th anniversary of his effect.

mechanics. In 1983, on the occasion of the 60th anniversary (Diamond Jubilee) of his effect, Hanle visited me at ETH Zurich (cf. Fig. 2). He was very excited that his effect was now, many decades after its discovery, being applied in astrophysics for the diagnostics of magnetic fields. The great value of the Hanle effect in astrophysics is that it provides information on magnetic fields in parameter domains, where the Zeeman effect is insensitive or even blind.

Nearly hundred years ago, the Zeeman effect was introduced to astrophysics for the first time by George Ellery Hale, who used it at Mt Wilson to discover magnetic fields in sunspots (Hale 1908). Since the introduction of the photoelectric magnetograph half a century ago (Babcock 1953), the Zeeman effect is regularly used to map solar magnetic fields and to study their evolution.

One of the giants in theoretical solar physics during the past several decades has been Gene Parker, who among many other things explained the supersonic solar wind and the principles of the solar dynamo. Figure 3 shows him at a COSPAR meeting in Tokyo in 1968 together with several other key persons in solar magnetic field work, like Bob Howard and John Wilcox, who initiated and developed fundamental synoptic programs for the Sun's magnetic field.



Figure 3. Gene Parker and fellow solar magnetic field fans in Tokyo, 1968.

2. Evolution of the fields

Macroscopic magnetic fields, on scales ranging from galaxies to planets, are believed to be produced by dynamo processes. The Sun is a prototypical dynamo that can be explored in detail. Like the galactic and planetary dynamos, the solar dynamo embodies the interaction between magnetic fields, turbulent convection, and rotation. Convection enhances the diffusion of the magnetic field lines by many orders of magnitude and transports angular momentum, which makes the rotation differential. The rotation breaks the left-right symmetry of the turbulence through the Coriolis force and makes the convective motions cyclonic, with opposite helicities in the two hemispheres. The differential rotation winds up the frozen-in field lines of an initial poloidal field to create a toroidal field. The cyclonic turbulence twists this toroidal field to regenerate a poloidal field. Through the turbulent diffusion, field lines get transported globally from local sources.

The resulting dynamo-produced magnetic field represents an oscillation between a poloidal and a toroidal state. The oscillation period is 22 yr (the Hale cycle), with sign reversal of the field every 11 yr. The sunspots represent portions of the toroidal field that have been carried up to the surface by the buoyancy forces. Therefore they appear as bipolar pairs oriented in the E–W direction, which have opposite polarity orientations in the two hemispheres. Each new 11-yr cycle the polarity orientations are reversed. These polarity rules are called Hale’s polarity law, which finds a natural explanation in terms of the mentioned dynamo scenario.

The evolving polar fields reach a maximum flux around the time of minimum sunspot activity, and represent a global N–S magnetic dipole with opposite polarities of the N and S polar caps. The polarity is the same as the polarity of the leading (W) portion of

bipolar sunspot groups in the same hemisphere. The polar fields reverse sign around the time of maximum sunspot activity. This reversal is, however, not synchronized between the N and S poles, a time difference of up to 2 yr is common.

As the Sun rotates with a period of approximately 27 days, we can build synoptic maps from the longitude band around the central meridian of each daily magnetogram to obtain the field distribution for all latitudes and longitudes. Such maps show a polarity pattern that is dominated by the magnetic fluxes in the sunspot latitude belts, with young, compact bipolar regions mixed with older, more dispersed bipolar flux that has been distorted by differential rotation. Due to the limited spatial resolution of such representations, however, very compact bipolar regions with much wider latitude distributions, like the numerous ephemeral active regions or the still smaller and more numerous intranetwork regions, are filtered out by the spatial smearing window.

Similar to the modal analysis of helioseismology, one way of exploring the global evolution of solar magnetic fields is to decompose the observed pattern in the synoptic maps into spherical harmonics, and then do time series analysis of the harmonic coefficients (Stenflo & Vogel 1986; Stenflo & Güdel 1988). The dominating property revealed by this analysis is a strict parity selection rule: The axially symmetric modes (with $m = 0$) are governed by the fundamental 22-yr period only when the parity is odd (the spherical harmonic degree l is odd, which means that the magnetic pattern is anti-symmetric with respect to reflections in the equatorial plane). The 22-yr period is practically absent in the even modes. Dynamo theories allow both odd and even modes to be excited, and it is not clear why the Sun has settled for such a strict choice of parity.

Although the general dynamo framework is not in serious doubt, many unanswered questions remain, mainly because of the complexity in trying to model magnetoconvection inside the Sun. The currently prevailing view is that most of the dynamo action takes place at the bottom of the convection zone, where the flux can be stored long enough before floating up to the surface. It is not clear how the Sun rids itself of all the magnetic flux that is built up each 11-yr cycle. Statistically the removal must be done as fast as the production rate.

3. Small-scale structuring

Magnetoconvection produces structuring over a dynamic range that covers many, possibly ten, orders of magnitude, from the global scales that govern the large-scale evolution of the 22-yr magnetic cycle, to the diffusion scales where the fields are destroyed. While we can resolve about four orders of magnitude of this scale spectrum, from 10^6 to 100 km, most of the structuring exists in the regime that will remain spatially unresolved in the foreseeable future. Fundamental physical processes occur at these unresolved scales, related to the emergence and destruction of flux, the heating processes and wave generation, and the fundamental turbulent processes leading to the fractal appearance of the structuring. Therefore spatially resolved observations always need to be combined with indirect diagnostics that provide information about the unresolved domain, something that is unfortunately often ignored.

The boundary between the resolved and unresolved domains gradually shifts towards smaller scales as advances in the resolving power of the instruments are made. Four decades ago, when I did the observations for my PhD with the magnetograph at

the Crimean Astrophysical Observatory in 1965–66, I noticed how the field strengths went up when the spectrograph slit size used for the observations was decreased (in height). The question therefore arose what the field strengths would be if we would have infinitely high resolution. This question of course could not be answered by direct observations, so the line-ratio technique was developed, which showed that more than 90% of the net quiet-Sun magnetic flux through the resolution elements had its origin in magnetic fields with strengths of 1–2 kG, although the apparent field strengths (as averaged over the resolution element) were only a few G (cf. Stenflo 1973, 1994; Stenflo *et al.* 1984). This result implied highly intermittent magnetic fields with tiny filling factors, which led to a whole industry of theoretical modelling in terms of narrow flux tubes embedded in a largely field-free atmosphere.

While the line-ratio technique did not provide direct information on the flux tube size, comparison with the number densities of spicules and bright points suggested that the flux tube diameters would be of order 100–300 km (Stenflo 1973). At that time this was far below the resolution limit of the existing telescopes. Several decades later the observational techniques have now advanced such that, with the help of image restoration and adaptive optics, we are beginning to actually resolve these kG flux tubes, which are being considered as basic building blocks of solar magnetism. Theoretical concepts of convective collapse of flux tubes (Parker 1978; Spruit 1979; Spruit & Zweibel 1979) have provided an explanation for the kG field strengths. Various observations, in particular with infrared lines (Solanki *et al.* 1996) have given support to the convective collapse mechanism while also showing the existence of a family of weaker flux tubes that had been predicted by Venkatakrishnan (1986).

Still the idealized flux tube scenario has been unsatisfactory, since it suggests that close to 99% of the photospheric volume (outside the kG flux tubes) would be field-free, which is nonsensical, since in the highly conducting and turbulent solar plasma magnetic fields must be ubiquitous. The circumstance that the magnetograms look empty between the kG flux fragments must either mean (i) that the field in the apparently empty regions is weaker than the noise threshold in the magnetograms, or (ii) that the field strength is above that threshold, but the field polarities are mixed on scales smaller than the spatial resolution used. By using the Hanle effect we now know that alternative (ii) is the correct one, and that the “hidden” field in the seemingly empty regions is actually relatively strong, of order 30–60 G throughout the volume (in contrast to the flux tube 1–2 kG occupying on average about 1% of the volume).

The Hanle effect is a coherency phenomenon that only occurs when coherent scattering contributes to the line formation. It represents all magnetic-field induced modifications of the scattering polarization. It has different field sensitivity and symmetry properties than the ordinary Zeeman-effect polarization. Thus the Hanle effect is sensitive to weaker fields, to horizontal fields in the solar chromosphere, and to small-scale fields with randomly mixed polarities within the spatial resolution element. The Hanle and Zeeman effects therefore complement each other. In particular, the Hanle effect can reveal the spatially unresolved “hidden” turbulent flux between the kG flux tubes (Stenflo 1982; Faurobert-Scholl 1993; Faurobert-Scholl *et al.* 1995; Stenflo *et al.* 1998), while the Zeeman effect is “blind” to these fields. The Stokes vector images in Fig. 4 show examples of Hanle signatures.

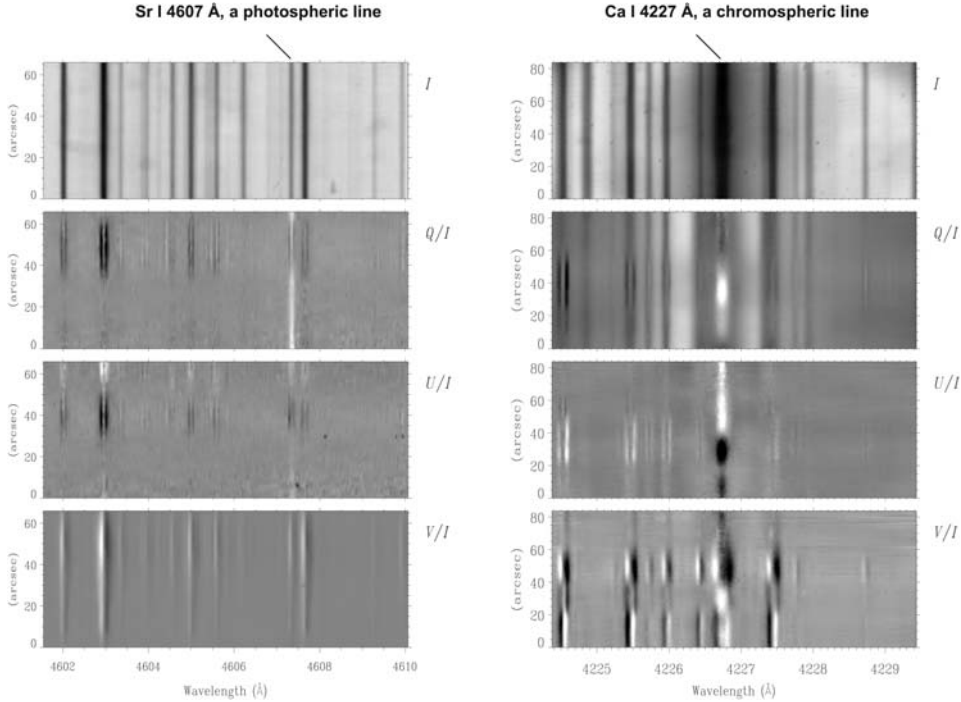


Figure 4. Recordings of the Stokes vector (the intensity I and the three fractional polarizations Q/I , U/I , and V/I) illustrating polarization signatures of the Hanle effect. The photospheric Sr I 4607 Å line and the chromospheric Ca I 4227 Å line exhibit strong linear polarization due to coherent scattering, which in the presence of magnetic fields gets modified by the Hanle effect (seen as spatial variations of Q/I and U/I along the slit). The surrounding spectral lines display the usual transverse Zeeman effect in the linear polarization (Stokes Q/I and U/I), the longitudinal Zeeman effect in the circular polarization (Stokes V/I). The recordings were made with the Zurich Imaging Polarimeter (ZIMPOL, cf. Povel 1995; Gandorfer *et al.* 2005) at the McMath-Pierce facility (Kitt Peak).

When looking at the world with our diagnostic tools, we see filtered versions of reality. When we put on our “Zeeman goggles”, we see the contributions from the kG flux tubes, but filter out the turbulent fields in between. When on the other hand we put on our “Hanle goggles”, we see the turbulent fields but not the kG flux tubes. We, however, increasingly realize that these two filtered versions of reality are merely two essential aspects of solar magnetism. A more unified, comprehensive view leads us to a reality that is fractal-like (Stenflo & Holzreuter 2003; Stenflo 2004). This view is supported by numerical simulations of magnetoconvection (Janssen *et al.* 2003). The pattern of quiet-Sun magnetic fields appears to maintain a high degree of self-similarity as we zoom in on ever smaller scales, as illustrated in Fig. 5. Instead of being limited to the idealized dichotomy of flux tubes and turbulent fields we need to work with probability distribution functions (PDFs) like in magnetoconvective theory. Application of PDFs to the interpretation of Hanle effect observations shows that the apparently field-free regions in solar magnetograms are filled with turbulent magnetic fields that carry magnetic energy that is significant for the overall energy balance of the solar atmosphere (Trujillo Bueno *et al.* 2004).

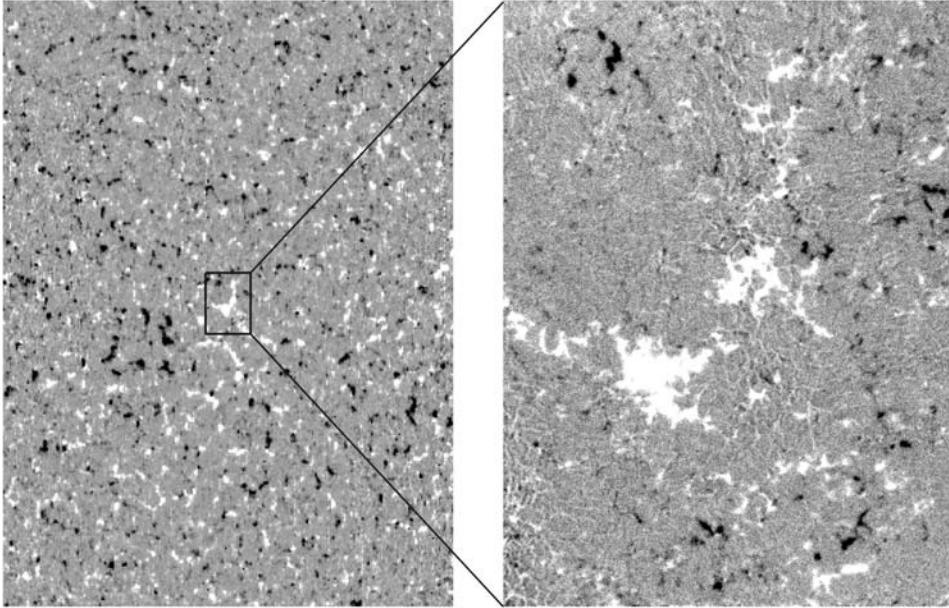


Figure 5. Illustration of the fractal-like pattern of magnetic fields on the quiet Sun. The rectangular area covered by the left map (from Kitt Peak) is about 15% of the area of the solar disk, while the map to the right (from the Swedish La Palma telescope) covers an area that is 100 times smaller. Bright and dark areas correspond to magnetic flux of positive and negative polarities, separated by grey voids of seemingly no flux. Analysis of Hanle-effect observations at IRSOL (Istituto Solari Locarno) of atomic and molecular lines have shown that these grey regions are actually no voids at all, but are teeming with turbulent magnetic fields.

4. Outlook

Synoptic programs are indispensable for learning about the global properties of the Sun. The program of daily magnetograms pioneered at Mount Wilson and extended at Kitt Peak has given us a unique dataset that now spans several solar cycles. From this we have learnt how the global evolution of the magnetic field pattern exhibits remarkable symmetries that constrain the nature of the dynamo. A new, significantly extended synoptic program, SOLIS, is now being implemented by the National Solar Observatory (USA), which will add a number of new aspects to the continued dataset, in particular vector magnetic field maps. A very important complement to the synoptic magnetic data comes from the MDI instrument on SOHO, which has provided us with the best-quality sequence of full-disk magnetograms for more than a solar cycle now.

Observations with improved angular resolution in combination with indirect diagnostics have shown that the magnetic structuring extends from the global scales towards the diffusion scales far below the telescope resolution. There is a remarkable degree of self-similarity over the various scales, indicating a fractal-like nature. A key to physical understanding may lie in the scales we are just beginning to resolve (the dimensions of the photon mean free path and the pressure scale height). The diagnostic methods to explore the unresolved scales are being extended, in particular through applications

of the Hanle effect. The interpretation of these indirect diagnostic data make use of idealized, parameterized models. The task of the interpretation is to determine the free parameters of the model (which for instance could describe the analytical form of the PDF distribution function). The choice of such interpretative models is guided by numerical simulations of magnetoconvection. Although such numerical simulations provide important guidance, they in turn need to be guided by observations, since the dynamic range that they can cover in any foreseeable future is only a small fraction of the full dynamic range of scales (possibly something like ten orders of magnitude) that are part of the Sun's magnetoconvection.

We may expect dramatic observational advances in the coming years. Some of the most burning questions concern the small-scale evolution of the magnetic field. The Hinode satellite offers unprecedented possibilities to explore this small-scale evolution and its relation to the thermodynamics. New ground-based facilities (SST, GREGOR, ATST, etc.) will allow resolution of the critical scales of the photon mean free path and the pressure scale height. It is important to combine these advances of high-resolution direct observations with the new opportunities that are now opening up to explore the spatially unresolved scales via the Hanle effect. Another new development is that coronal magnetic fields are becoming measurable (from ground and from space). The interpretation of all these new observations will be greatly aided by the significant advances that can be expected in the numerical simulation of magnetoconvection and other MHD processes.

Let me end by summarizing what I consider to be some of the most important observational questions that we need to answer to understand the operation of the solar dynamo. They concern the emergence, decay, and removal of the photospheric magnetic flux. In particular: What is the fate of all the emerged magnetic flux? It has to be removed on the solar cycle time-scale and statistically be in equilibrium with the emergence rate, otherwise the photosphere would quickly be choked with undisposed magnetic flux. We know that as we go from active regions to ephemeral regions and still smaller scales, the emergence rate increases dramatically, but it is not so easy to see how one can get rid of the flux at such tremendous rates. Which is the main process? Basically we need to determine the relative contributions of the following three alternatives: (i) Cancellation of opposite polarities (reconnection). (ii) Flux retraction (reprocessing in the convection zone). (iii) Flux expulsion (and the role of CMEs).

In spite of decades of hard work we still know practically nothing about the relative roles of these processes. With the new observational tools, however, we are on the verge of significant progress in this fundamental area.

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Global Solar Dynamo Models: Simulations and Predictions

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Abstract. Flux-transport type solar dynamos have achieved considerable success in correctly simulating many solar cycle features, and are now being used for prediction of solar cycle timing and amplitude. We first define flux-transport dynamos and demonstrate how they work. The essential added ingredient in this class of models is meridional circulation, which governs the dynamo period and also plays a crucial role in determining the Sun's memory about its past magnetic fields. We show that flux-transport dynamo models can explain many key features of solar cycles. Then we show that a predictive tool can be built from this class of dynamo that can be used to predict mean solar cycle features by assimilating magnetic field data from previous cycles.

Key words. Sun—magnetic fields: solar cycle—dynamo—prediction.

1. Introduction and philosophical approach

In the current era of extensive use of the high atmosphere and neighboring interplanetary medium by man, solar cycle predictions have considerable practical value. Starting with solar cycle 22 (1986–96), NASA and other Federal Agencies concerned with the effects of solar activity on human activity have convened cycle prediction panels to assess all predictions and methods used, and give their best judgment on the amplitude and other properties of the next cycle. For cycles 22 and 23, prediction methods were primarily statistical rather than dynamical. That is, no physical laws were integrated forward in time, as is done for meteorological and climate predictions. On the verge of the onset of solar cycle 24, the first such ‘dynamo based’ cycle prediction, which involves integrating forward in time a form of Faraday’s law of electromagnetic theory, has now been made (Dikpati *et al.* 2006; Dikpati & Gilman 2006).

Given the complexity of the physics in the interior of the Sun, how is such a prediction possible? Any such model must be able to simulate the important features of a ‘typical’ solar cycle (Dikpati *et al.* 2004), such as the ‘butterfly diagram’, and the phase relation in time between toroidal (sunspot) fields and poloidal (polar) fields. It must do this with differential rotation and meridional circulation that are close to that observed in the Sun. In other words, the dynamo model must be ‘calibrated’ with solar observations. This process also constrains the choice of more poorly known physical parameters, such as the magnetic diffusivity. Once calibrated, the model can be used to simulate and predict

departures from the typical cycle, caused by variations in various input data, such as observed magnetic fields from previous times. We and others have therefore focussed attention on a class of dynamo model that has enough physics in it to be calibratable with global solar observations, while at the same time being integratable using modest computing resources. This class of dynamo is the ‘flux transport’ dynamo, so named (Wang & Sheeley 1991) because of the inclusion of meridional circulation as well as differential rotation in the induction equation. Such models generally parameterize key turbulent processes occurring on scales smaller than the resolution of the model, such as the turbulent transport of magnetic flux captured by a turbulent magnetic diffusivity, and the turbulent lifting and twisting of magnetic field lines represented by one or more ‘ α -effects’. These models also do not include the buoyant eruption of toroidal flux tubes, turbulent diamagnetic effects and Alfvén wave effects. The success of such a model in simulating and predicting solar cycles can only be judged by studying the output from it, by comparing it to the appropriate solar observations. Figure 1 shows a sequence of schematic diagrams that depict qualitatively the succession of processes contained in a flux-transport dynamo solution that leads to cyclic magnetic fields of the correct dominant symmetry about the equator. The historical evolution of large-scale solar dynamo models, starting from the first model by Parker (1955) was reviewed by Dikpati (2004b). Here we focus primarily on flux-transport dynamo mechanism and the predictive tool development from it.

2. Mathematical formulation

We start from the axisymmetric, mean-field dynamo equations and write them as:

$$\begin{aligned} \frac{\partial A}{\partial t} + \frac{1}{r \sin \theta} (\mathbf{u} \cdot \nabla) (r \sin \theta A) \\ = \eta \left(\nabla^2 - \frac{1}{r^2 \sin^2 \theta} \right) A + S(r, \theta, B_\phi) + \frac{\alpha B_\phi}{1 + (B_\phi/B_0)^2}, \end{aligned} \quad (1)$$

$$\begin{aligned} \frac{\partial B_\phi}{\partial t} + \frac{1}{r} \left[\frac{\partial}{\partial r} (r u_r B_\phi) + \frac{\partial}{\partial \theta} (u_\theta B_\phi) \right] \\ = r \sin \theta (\mathbf{B}_p \cdot \nabla) \Omega - \hat{\mathbf{e}}_\phi \cdot [\nabla \eta \times \nabla \times B_\phi \hat{\mathbf{e}}_\phi] + \eta \left(\nabla^2 - \frac{1}{r^2 \sin^2 \theta} \right) B_\phi, \end{aligned} \quad (2)$$

in which $B_\phi(r, \theta, t) \hat{\mathbf{e}}_\phi$ and $\nabla \times A(r, \theta, t) \hat{\mathbf{e}}_\phi$ respectively represent the toroidal and poloidal fields.

In equation (1), $S(r, \theta, B_\phi)$ represents the production of new poloidal field from the decay of tilted, bipolar active regions at the surface, and α is an α -effect that arises from global instabilities in the overshoot tachocline (see Dikpati & Gilman 2001 for details). In equation (2), $\Omega(r, \theta, t)$ is the specified differential rotation, parameterized from helioseismology, and in both equations $\mathbf{u} = u_r \hat{\mathbf{e}}_r + u_\theta \hat{\mathbf{e}}_\theta$ is the specified meridional circulation. In equation (1), α now is an α -effect that arises from instabilities in the overshoot tachocline near the bottom of the domain. The details of the tachocline α -effect we have used, such as its latitudinal profile, are given in Dikpati & Gilman (2001). Both the surface poloidal source and the α -effect near the bottom of the domain are ‘quenched’ above a certain field strength, so that the

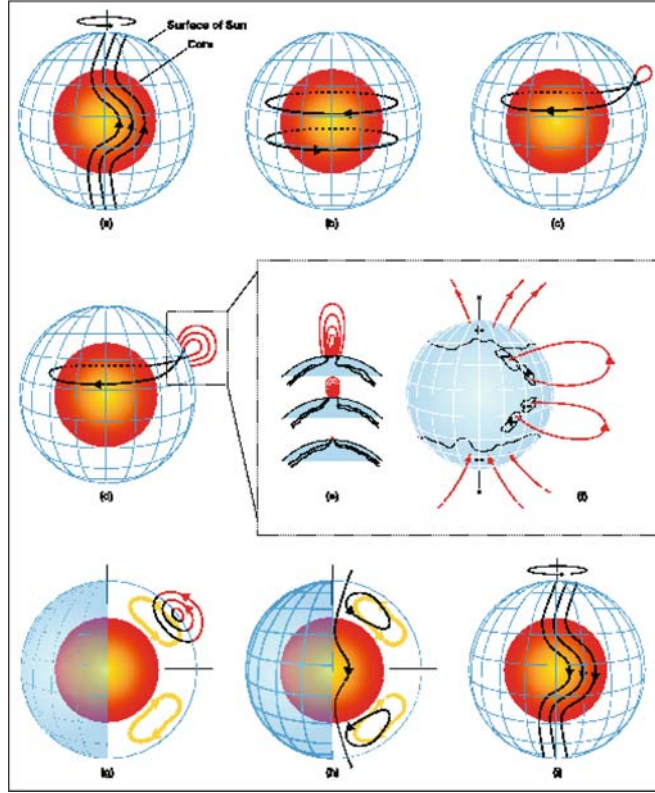


Figure 1. Schematic of solar flux-transport dynamo processes. Red inner sphere represents the Sun's radiative core and blue mesh the solar surface. In between is the solar convection zone where dynamo resides. (a) Shearing of poloidal field by the Sun's differential rotation near convection zone bottom. The Sun rotates faster at the equator than the pole. (b) Toroidal field produced due to this shearing by differential rotation. (c) When toroidal field is strong enough, buoyant loops rise to the surface, twisting as they rise due to rotational influence. Sunspots (two black dots) are formed from these loops. (d, e, f) Additional flux emerges (d, e) and spreads (f) in latitude and longitude from decaying spots (as described in Fig. 5 of Babcock 1961). (g) Meridional flow (yellow circulation with arrows) carries surface magnetic flux poleward, causing polar fields to reverse. (h) Some of this flux is then transported downward to the bottom and towards the equator. These poloidal fields have sign opposite to those at the beginning of the sequence, in frame (a). (i) This reversed poloidal flux is then sheared again near the bottom by the differential rotation to produce the new toroidal field opposite in sign to that shown in (b).

solutions do not grow without bound. This is the only nonlinearity in the system. The boundary conditions we apply are: the toroidal field is zero on all boundaries, while the poloidal field is fitted to a potential field above the top boundary. $A = 0$ at the poles is required by axisymmetry. $A = 0$ at the bottom boundary ($r = 0.6R$) keeps the field out of the deep interior, and in the case of a single hemisphere calculation, $B_\phi = 0$; $\partial A / \partial \theta = 0$ at the equator to ensure the antisymmetric magnetic field solutions. In full sphere calculations, the equatorial boundary conditions are omitted.

3. Major results since 1999

3.1 *What determines the period of the solar activity cycle?*

A principal result of Dikpati and Charbonneau (1999) was to show that in flux transport dynamos, the dynamo period is approximately inversely proportional to the meridional flow speed. Agreement between the model cycle period and that of the Sun requires an equatorward return flow of a few m/s near the base of the convection zone. The cycle period in flux-transport dynamos is rather insensitive to the amplitude of the differential rotation, α effects, or the diffusivity, in contrast to that in the so-called interface dynamos (Parker 1993) and classical α - ω dynamo (Stix 1973).

3.2 *What determines the phase relation between the tachocline toroidal field and the surface radial field?*

Wang *et al.* (1989) first showed, using a one-dimensional flux-transport model applied to the Sun's surface, that advective-diffusive transport of the Sun's weak magnetic flux arising from the decay of active regions can explain quantitatively the poleward drifts of these flux and the polar reversal during sunspot maximum. This was confirmed by simulations from a complementary 2D flux-transport model by Dikpati (1996) for a meridional cut in the Sun's convective envelope. In a detailed Babcock-Leighton flux-transport dynamo calculation with solar-like differential rotation, Dikpati and Charbonneau (1999) showed that this phase relation is a natural consequence of the meridional flow of observed amplitude on the Sun. The key to getting the correct phase relation between the peak in the toroidal fields and sunspots, and the surface poloidal fields, particularly the polar field reversals is having the primary creation of toroidal fields occurring near $r = 0.7R$, while the poloidal flux is created at the surface and transported towards the pole by surface processes.

3.3 *What causes the solar dynamo to operate in a mode predominantly antisymmetric about the equator, as exemplified by Hale's polarity laws?*

Babcock-Leighton type advection-dominated flux-transport dynamos cannot produce solar-like antisymmetric magnetic fields about the equator. By doing full spherical shell simulations without imposing equatorial symmetry conditions, Dikpati & Gilman (2001) showed that the combination of a meridional flow towards the equator at the bottom of the convection zone, together with an α -effect there, favors the antisymmetric mode.

The physical reason why an advection-dominated Babcock-Leighton dynamo does not relax to an antisymmetric mode is that the poloidal fields generated at the surface diffuse away in the course of their long traversal to meet with their opposite-hemisphere counterparts and to form an extended equatorial dipole, which is necessary for producing antisymmetric toroidal fields. In contrast, poloidal fields produced by a tachocline α -effect are swept towards the equator quickly by the equatorward meridional flow there, and hence can couple the two hemispheres by antisymmetric magnetic fields.

3.4 *Can flux-transport dynamos explain features of global solar magnetism specific to a particular solar cycle?*

Dikpati *et al.* (2004) was the first to apply a flux-transport dynamo to explain specific features of a particular solar cycle, that of cycle 23. The model was first calibrated against an average solar cycle, and then initialized with magnetic data from cycle 22. They found that the 1.5-year delay in polar field reversal in cycle 23 compared to an average cycle was primarily due to the weaker surface poloidal field sources in cycle 23 compared to 22. When transported to the poles by the meridional circulation, these weaker sources took longer to cancel out the polar field created during cycle 22. These weaker sources also caused the post-reversal polar field of cycle 23 to be weaker than in cycle 22, and take longer to build up. Dikpati *et al.* (2004) also explained that the $\sim 50\%$ slow-down in the meridional flow-speed during 1996–2003 (Haber *et al.* 2002; Basu & Antia 2003) together with the N-S asymmetry in this flow, caused the North Pole reverse ~ 1 year before the South Pole.

In response to Choudhuri and collaborators’ “million-dollar question” raised in various emails spread over the world (including in the Diamond Jubilee conference at PRL) about the correctness of Dikpati *et al.* (2004) results, we (Dikpati & Gilman 2007) recently published the steps for building such a model (see Fig. 2) and re-confirmed the results of the calibrated dynamo of Dikpati *et al.* (2004).

3.5 *Can the so-called interface dynamos calibrate well to the Sun?*

Parker (1993) defined and applied the first interface dynamo to the Sun. In its simplest form this dynamo contains a radial differential rotation characteristic of the solar tachocline, together with an α -effect just above the tachocline. The magnetic diffusivity is taken to be much smaller in the tachocline than above it. Parker showed that it is capable of producing large amplitude toroidal fields in the tachocline due to the low diffusivity, where they can be stored long enough to reach the strength needed to be the source of sunspots seen in middle and low latitudes.

But does such a solution calibrate well with solar observations? For example, does it produce a butterfly diagram from toroidal field near $r = 0.7R$ and surface radial field that agrees well with that of the real Sun? Also, Garaud (1999) showed that the skin-depth of cyclically reversing poloidal fields near the base of the convection zone can be severely limited by the rapidly declining magnetic diffusivity there. What if conditions are such that these poloidal fields can not penetrate the tachocline enough to be sheared by the radial rotation gradient there? In that case, a pure interface dynamo might fail completely. Dikpati *et al.* (2005) showed in the absence of tachocline radial shear participating in the dynamo process, a latitudinal differential rotation can provide the necessary Ω -effect to drive an oscillation in an interface dynamo, but it alone cannot produce the latitudinal migration. They concluded that a flux-transport dynamo driven by both the Babcock-Leighton and interface/bottom α -effects is a robust dynamo for the Sun.

3.6 *Can the currently operating solar dynamo be a significant source of toroidal flux in the solar interior below the convection zone?*

One of the remaining mysteries concerning radiative stellar interiors is whether they contain magnetic fields, and if so, what kind and of what magnitude. Two concepts are

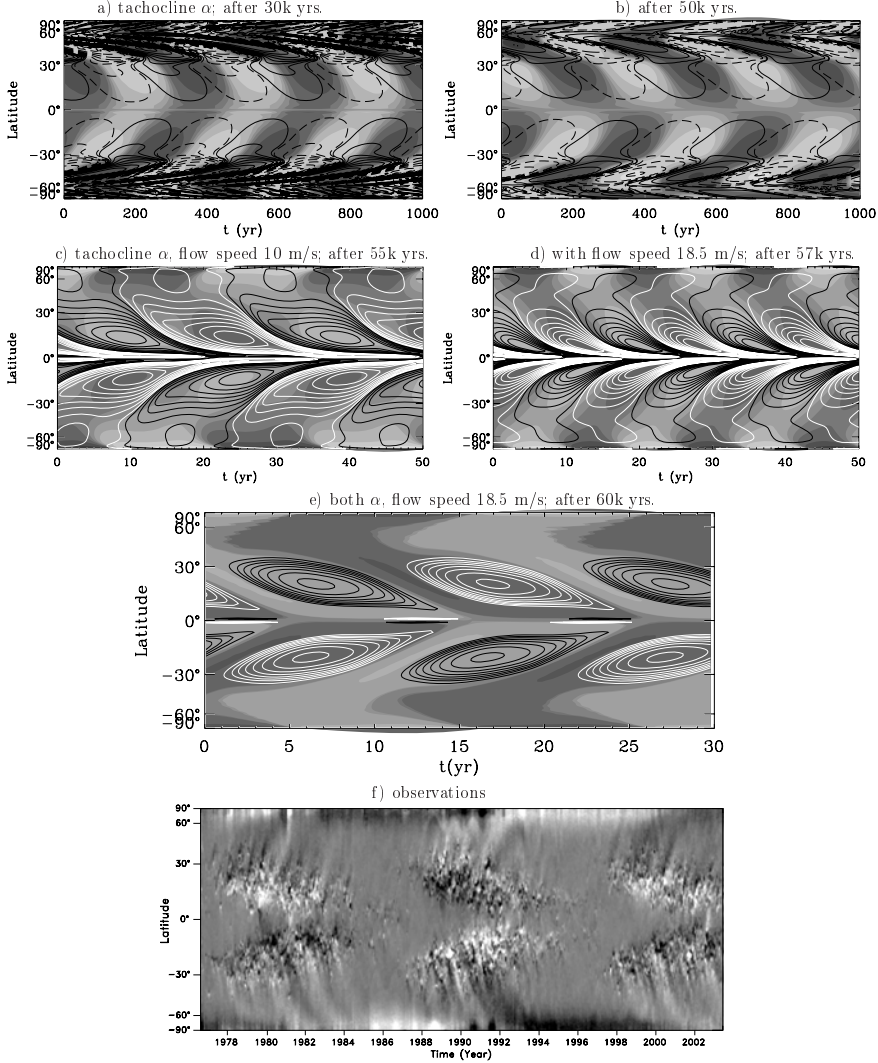


Figure 2. (a, b): Butterfly diagrams for $B_r|_R$ (gray shades) and $B_\phi|_{ov}$ (contours) after 30 k (less evolved), 50 k years, with differential rotation, tachocline α -effect but no meridional circulation. (c, d): Butterfly diagrams with meridional circulation (u_{mc}) added; (c) for maximum $u_{mc} = 10 \text{ ms}^{-1}$ after another 5 k years' run; (d) maximum $u_{mc} = 18.5 \text{ ms}^{-1}$, after additional 2 k years' run. Here the cycle period is governed by meridional circulation. (e): Butterfly diagram with Babcock-Leighton α -effect added, after 3 k years' run (though 50 years is enough). $B_\phi|_{ov}$ and $B_r|_R$ patterns in latitude show very good agreement with surface observations (f) (from Fig. 8 of DDGAW).

that such fields are primordial (Cowling 1945; Wrubel 1952) or that they are remnant fields from extinct dynamos (Schüssler 1975; Parker 1981). Mestel and Weiss (1987) suggested that convection zone dynamos with cycles fluctuating in amplitude could create net interior fields. Dikpati *et al.* (2006) show that flux-transport dynamos producing very regular cycles in the convection zone can create substantial non-reversing

toroidal fields in the outer part of the stellar interior, provided there is an α -effect near the bottom of the convection zone. Their results should apply to stellar interiors generally.

3.7 Can the timing of the end of a cycle and the onset of the next one be predicted?

Since for flux-transport dynamos the period is set primarily by the amplitude of the meridional flow, if that flow changes from one cycle to the next or within a given cycle, the length of that cycle will be affected. Starting from Dikpati *et al.*'s (2004) calibrated dynamo, Dikpati (2004a) performed a series of numerical experiments starting from cycle 22 and allowing for different meridional flow amplitudes during cycle 23. The resulting time trace of the induced toroidal field in the model tachocline for the different meridional flows is shown in Fig. 3. Here we see that, compared to a meridional flow set by the average cycle period for cycles 12–22 (10.75 yrs), if the meridional flow decelerates during 1996–2003 and then remains low, the onset of cycle 24 will be delayed by about 12 months, leading to new cycle spots exceeding the old cycle spots in early 2008. If the meridional flow instead accelerates after 2003, some of the delay can be made up, resulting in a net delay of about 6 months. This prediction of late onset of cycle 24 is now supported by a variety of observational evidence. For example, the observational butterfly diagram from NASA shown in Fig. 3(c) shows clearly that we have not yet (at the start of 2007) reached minimum. There is still significant cycle 23 activity in the Southern Hemisphere. In addition, the corona is not yet in minimum

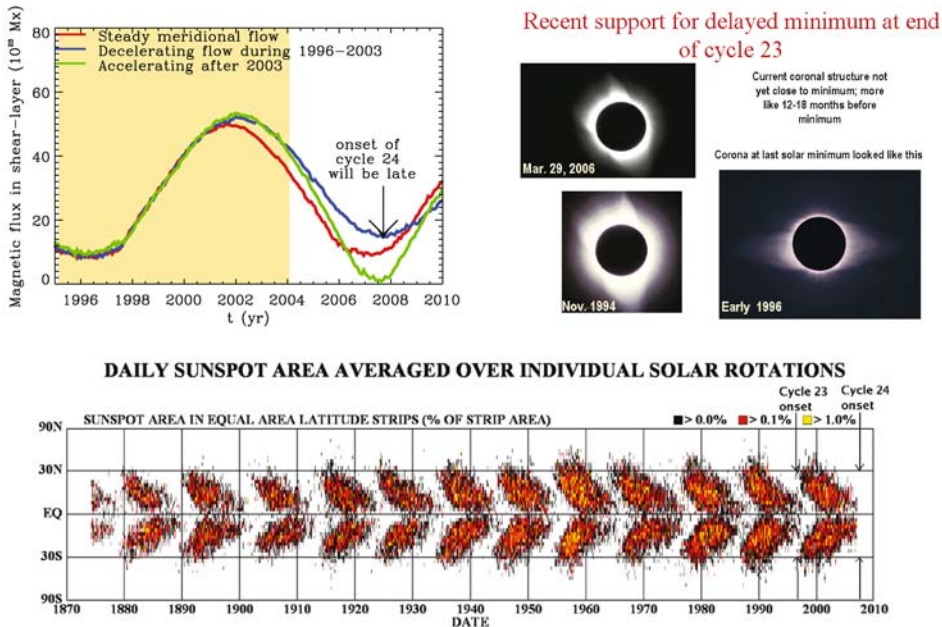


Figure 3. (a) Time trace of integrated toroidal flux in bottom shear layer for simulations beginning in 1995 with 3 different meridional flow variations with time; (b) Coronal images at eclipse, showing 2006 corona not at minimum phase; (c) Observed butterfly diagram from David Hathaway's website (<http://solarscience.msfc.nasa.gov/imagesbfly.gif>).

phase. This we can see in Fig. 3(b), which shows that the corona at total eclipse of March 29, 2006 looked very much like the corona of November 1994, not the dipolar corona of early 1996, during the last minimum. As the onset of a new cycle occurs ~ 1 year after the minimum, both of these observations are consistent with the next solar cycle onset being delayed until late 2007 or early 2008.

3.8 Can the sequence of observed solar cycle peaks be simulated and predicted?

The first truly dynamo-based simulation of the sequence of peaks of past solar cycles and a forecast of solar cycle 24 was published by Dikpati & colleagues (Dikpati *et al.* 2006; Dikpati & Gilman 2006). These simulations and forecast were done by starting from the calibrated flux transport dynamo of Dikpati *et al.* (2004) and making certain modifications. First, in equation (1) for the poloidal field in terms of vector potential A , the link between the toroidal field in the dynamo at the bottom and the poloidal field at the surface is replaced by a surface forcing term that is derived from the waxing and waning of the observed poloidal fields created by the decay of tilted bipolar active regions in previous solar cycles. This represents a form of 2D (time-latitude) data assimilation in the model, a very common practice for meteorological forecast models, but perhaps the first time it has been used in a solar forecasting problem.

As a result of the above modification, no quenching of the Babcock-Leighton poloidal source, and no buoyancy mechanism are included in the model. However, the nonlinear quenching in the α -effect near the base of the convection zone is retained. Since this α -effect is much smaller compared to the Babcock-Leighton surface α -effect, in effect the model is mathematically changed from being a self-contained nonlinear system to a primarily linear system that is forced at the upper boundary by a measure of the Sun's past surface fields. The resulting equation for A is given by

$$\begin{aligned} \frac{\partial A}{\partial t} + \frac{1}{r \sin \theta} (\mathbf{u} \cdot \nabla) (r \sin \theta A) \\ = \eta \left(\nabla^2 - \frac{1}{r^2 \sin^2 \theta} \right) A + \mathcal{F}(r, \theta, t) + \frac{\alpha B_\phi}{1 + (B_\phi/B_0)^2}, \end{aligned} \quad (3)$$

in which B_0 is the quenching field strength, set to 30 kGauss.

In this class of models, primarily the meridional flow speed determines the cycle period. However, we do not have observations of the time variation of this flow prior to 1996. So, in our initial approach, we estimate an average amplitude of meridional flow from the average period (10.75 yrs) of as many past cycles as we have high quality sunspot data for, which takes us back to cycle 12. If we use a meridional flow amplitude determined in this way, but force the model with data that has the observed period of each sunspot cycle, a phase mismatch in the link between the toroidal and poloidal fields grows to an unacceptably large value over several cycles' time integration. This makes simulation and forecast of cycle peaks impossible. To avoid this problem, Dikpati, deToma and Gilman stretched and compressed in time the observed data so that each cycle in the surface forcing had the same period, 10.75 years, as the intrinsic period of the dynamo, to avoid destructive phase mismatches. Thus this forecast model can predict amplitude, but not length simultaneously. A more sophisticated approach might involve an ensemble of 12 cycle simulations with different or time-varying meridional circulation, but that remains for the future.

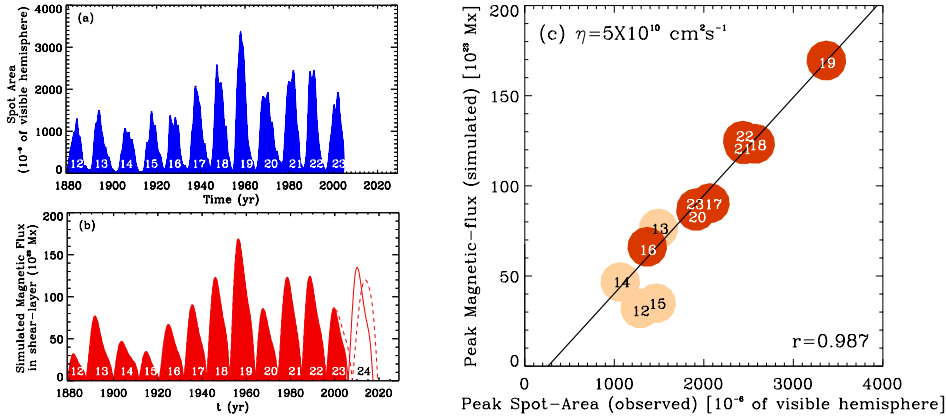


Figure 4. (a) Observed sunspot area data for cycles 12–23; (b) integral from 0–45° latitude of simulated toroidal magnetic flux in bottom shear layer for cycles 12–24, plus two forecasts for cycle 24; (c) scatterplot of sunspot area peaks vs. peak of toroidal magnetic flux integral.

Starting from a fully converged calibrated model solution, we initialized the model from the beginning of solar cycle 12, and ran it for the next 13 cycles extending through the upcoming cycle 24. The results of our simulations and predictions for the peaks of cycles 12–24 are shown in Figs. 4(a) and (b), adapted from Dikpati *et al.* (2006). Comparing the observed solar cycles with the sequence of simulated peaks in Fig. 4(a), we see that they are in good agreement, particularly beyond cycle 15 – the first few cycles the model needs to load its conveyor belt and to create its memory about the past cycles. The solid curve represents the case with meridional circulation fixed at a value that produces the 10.75 years dynamo period; it shows cycle 24 should have a $\sim 50\%$ higher peak than cycle 23. The dashed curve represents a simulation that takes account of time variations in observed meridional circulation since 1996, as described above in section 3.7. For this case, cycle 24 is forecast to be $\sim 30\%$ higher than cycle 23. There are many other forecasts for cycle 24, most saying it will be larger than 23, but some saying it will be smaller. Our forecast is the very first one made using a dynamo model with real solar data. The high skill of our forecast model is demonstrated in Fig. 3(b), which plots the observed peaks for cycles 12–23 against the peaks of the simulated toroidal flux in the overshoot tachocline between the equator and 45°. How the model achieves its forecast skill is discussed in detail in Dikpati and Gilman (2006). A key element is shown in Fig. 5 which plots the latitudinal poloidal field and the toroidal field at similar phases in cycles 19 and 20. It is the latitudinal fields from past 3 cycles, lined-up in the conveyor belt, that combine to form the seed for the next sunspot cycle’s spot-producing flux, in contrast to the polar fields from last minimum.

3.9 Criticisms and future calculations

Certain criticisms have been put forward concerning forecasting solar cycle peaks using flux-transport dynamo models by Tobias, Hughes and Weiss. The authors of this letter state *a priori* that because they contain parameterizations, flux-transport dynamos have no predictive power. But all weather and climate forecasting models

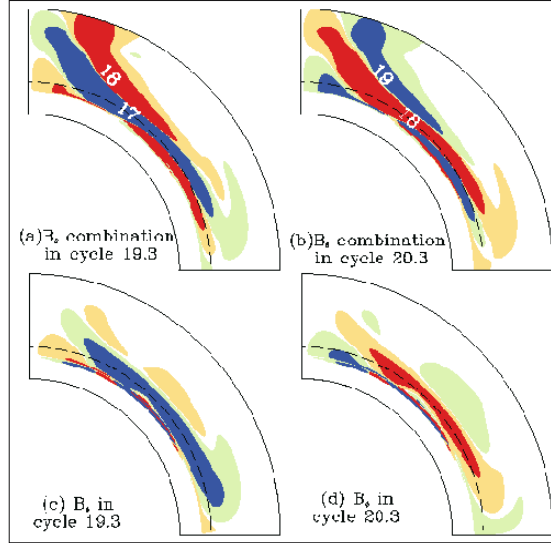


Figure 5. (a, b) Latitudinal component of simulated poloidal field (red and blue opposite signs) at phase 0.3 of cycles 19 and 20; (c, d) same for simulated toroidal field.

also have many parameterizations, and they have long been demonstrated to have skill. It is a logical certainty that predictive skill of ALL models, not just flux transport dynamos, can only be judged *a posteriori* by the results they produce. Also the skill of any model cannot be tested by using some other model with other physics, such as one that omits meridional circulation or time-latitude data assimilation. Much of the uncertainty of the effect of poorly known parameterizations of physical processes is removed by calibrating the model to solar observations, and then forecasting only departures from the calibrated solutions. This is a common and successful practice for forecast models in other fields, such as meteorology.

Tobias, Hughes and Weiss also claim that since the dynamo equations are extremely nonlinear, chaos intrinsic to the system makes prediction much more difficult, particularly for longer time periods. Within a solar cycle, there are short-term turbulent features which are probably governed by a variety of nonlinear processes including the buoyant rise of initially toroidal flux tubes. But we are not attempting to forecast them with our forced linear predictive tool. Obviously there are aggregate effects of many short-term nonlinear events which lead to, for example, the average surface poloidal fields that have a well-defined observed pattern. Our predictive tool captures this pattern. By analogy, mean flow in a river can be forecast without forecasting accurately the various eddies that occur in the flow.

Many additional calculations can and should be done with our solar cycle prediction model to test its robustness with respect to changes and fluctuations with time and space in various parameters currently included in it, such as the meridional circulation, turbulent magnetic diffusivity, α -effects, as well as the form and time dependence of the surface magnetic data from previous cycles. Such tests probably would be best done by employing so-called ‘ensemble-forecast’ methods, such as described in a meteorological context by Kalnay (2003), as well as more sophisticated methods

of assimilating observed data and iterating with updated forecast models, such as described in Kalnay (2003) and Talagrand (1997).

4. Concluding remarks

We have shown above that flux-transport dynamos when applied to the Sun can answer many important questions, from what determines the dynamo period to how the fields of previous cycles determine the amplitude of the next one. We plan many sensitivity tests of the model, and expect to generalize it, by using sequential data assimilation techniques to ultimately forecast period, amplitude and shape of a cycle. We already know that when we split the observed surface magnetic data into north and south hemispheres, we are able to correctly simulate the large differences in amplitude between hemispheres in later cycles when they occur.

Acknowledgements

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Prospects for Predicting Cycle 24

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Abstract. Although we have reliable data of solar polar fields only from the mid-1970s, it seems that the polar field at a minimum is well correlated with the next cycle, but the strength of the cycle is not correlated with the polar field produced at its end. We explain this by suggesting that the Babcock–Leighton mechanism of poloidal field generation from tilted active regions involves randomness, whereas the other aspects of the dynamo process are more ordered. To model actual cycles, we have to ‘correct’ our theoretical dynamo model by ‘feeding’ information about the polar field at the minima. Following this process, we find that our model fits the observed sunspot numbers of cycles 21–23 reasonably well and predicts that cycle 24 will be the weakest in a century.

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

1. Background

When we try to predict a solar cycle a few years ahead of time, have we reached the stage when we can regard our methodology to be slightly more reliable than betting in a horse race? That is the question we wish to discuss in this presentation.

Nearly three decades ago, it was suggested by Schatten *et al.* (1978) that the polar field at the preceding minimum gives an indication of the strength of the next solar cycle. If the polar field at a minimum is stronger than the usual, then the next cycle is likely to be strong, and *vice versa*. Let us discuss whether this method can be considered sufficiently reliable. We have systematic polar field data only from the mid-1970s when WSO and MWO started regular observations of the polar field. Svalgaard *et al.* (2005) have computed a quantity DM, which is a proxy of the dipole moment of the Sun and which is a good measure of the polar field of the Sun at the minimum. Table 1 lists DM values at the end of cycles 20–23, along with the strengths of cycles 20–23. In Fig. 1, we plot the strengths of the cycle $n + 1$ against the DM values at the end of cycles n . Although we have only 3 data points by now and one may justifiably question the statistical significance of any correlation based on 3 data points, all the 3 points lie so close to a straight line that it is difficult to believe that this is merely a statistical coincidence. If the solid straight line in Fig. 1 is assumed to be an approximate representation of a real correlation, then it is easy to predict the strength of the coming cycle 24. Since DM value at the present minimum is about 120 (indicated by the dashed line), the predicted strength of cycle 24 can be merely read off from the vertical scale corresponding to the point where the dashed line cuts the solid line

Table 1. Maximum strength of the cycle and the DM value at its end are listed against cycle number. The data for this table are taken from Svalgaard *et al.* (2005).

Cycle number	Maximum strength of the cycle $R_{\max}$	DM value at the end of cycle
20	110.6	250
21	164.5	245.1
22	158.5	200.8
23	120.8	119.3

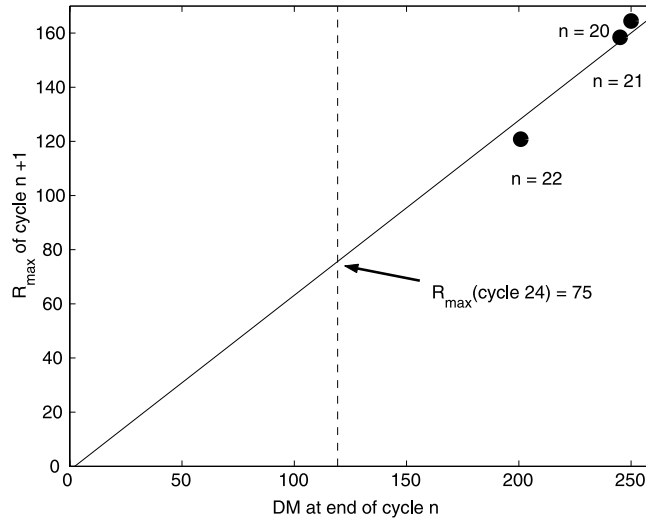


Figure 1. Strengths of solar cycles plotted against the DM values of polar fields at the preceding minima.

representing the correlation. This gives an $R_{\max}$ of about 75, suggesting that cycle 24 will be the weakest in a century. This is essentially the basis of the predictions made by Schatten (2005) and Svalgaard *et al.* (2005), who believe that cycle 24 will be very weak.

We now do the next interesting thing of plotting DM at the end of cycle n against the strength of that cycle. This is shown in Fig. 2. It is clear even from the limited number of data points that there is no good correlation. This means that the strength of the cycle does not completely determine the polar field at the end of the cycle. A strong cycle is not necessarily followed by a strong polar field at its end. Figure 1 virtually suggests that the polar field at the end of a cycle in some way almost determines the next cycle. But we conclude from Fig. 2 that the strength of the cycle does not determine the polar field at its end in the same causal way. There is a break or at least a weakening in the chain of causality when we go from the maximum of a cycle to the following minimum. We actually need such a break in strict causality somewhere in the cycle because the solar cycles are irregular. If strict causality were followed everywhere, we would have expected the cycles to be much more regular. We shall further discuss the implications of these figures later in our presentation.

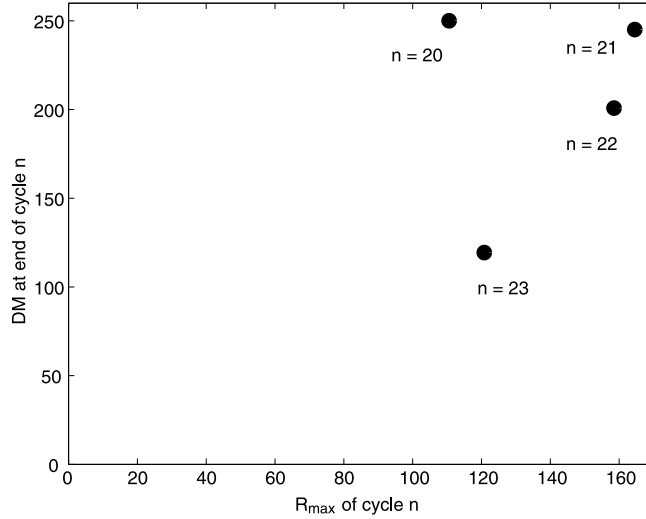


Figure 2. DM values of polar fields at the minima plotted against the strengths of the previous solar cycles.

Since the solar cycle is believed to be produced by a hydromagnetic dynamo process, we would now like to raise the question whether the solar dynamo theory can be applied to predict solar cycles and whether the empirical method of using the polar field strength at the preceding minimum as an indicator for the next cycle can be justified theoretically from dynamo models. Only within the last few years, solar dynamo models have become sufficiently sophisticated and realistic to address these issues. Dikpati & Gilman (2006) have published the first prediction of cycle 24 based on a solar dynamo model. They claim that cycle 24 will be the strongest in 50 years. If their prediction turns out to be true, then the fourth data point in Fig. 1 will lie on the dashed line a little above the upper boundary of the figure, far away from the solid line which indicates a possible correlation! Several dynamo theorists have been harshly critical of the work of Dikpati & Gilman (2006). Tobias *et al.* (2006) make the following comment on this work: “Any predictions made with such models should be treated with extreme caution (or perhaps disregarded), as they lack solid physical underpinnings.” While we are also highly critical of the Dikpati–Gilman work for reasons which we shall explain in section 3, we cannot also accept the opposite extreme viewpoint of Tobias *et al.* (2006), who suggest that the solar dynamo is a nonlinear chaotic system and predictions are impossible or useless. If that is the case, then we are left with no explanation for the correlation seen in Fig. 1.

This state of affairs made us feel that it was important to take a fresh look at the problem from a new angle. Our first results on this subject (Choudhuri *et al.* 2007) are discussed in section 2. Finally section 3 presents our concluding remarks, contrasting our methodology with that of Dikpati & Gilman (2006).

2. Our model

Some processes in nature can be predicted and some cannot. We can easily calculate the trajectory of a projectile by using elementary mechanics. On the other hand, when

a dice is thrown, we cannot predict which side of the dice will face upward when it falls. Is the solar dynamo more like the trajectory of a projectile or more like the throw of a dice? Our point of view is that the solar dynamo is not a simple unified process, but a complex combination of several processes, some of which are predictable and others not. Let us look at the processes which make up the solar dynamo.

Most of the current solar dynamo models combine three basic processes:

- (i) The strong toroidal field is produced by the stretching of the poloidal field by differential rotation in the tachocline.
- (ii) The toroidal field generated in the tachocline gives rise to active regions due to magnetic buoyancy and then the decay of tilted bipolar active regions produces the poloidal field by the Babcock–Leighton mechanism.
- (iii) The meridional circulation advects the poloidal field first to high latitudes and then down to the tachocline.

Two-dimensional mean field dynamo models based on these three processes were first constructed about a decade ago (Choudhuri *et al.* 1995; Durney 1995). We believe that the processes (i) and (iii) are reasonably ordered and deterministic. In contrast, the process (ii) involves an element of randomness due to the following reason. The poloidal field produced from the decay of a tilted bipolar region by the Babcock–Leighton process depends on the tilt. While the average tilt of bipolar regions at a certain latitude is given by Joy’s law, we observationally find quite a large scatter around this average. Presumably the action of the Coriolis force on the rising flux tubes gives rise to Joy’s law (D’Silva & Choudhuri 1993), whereas convective buffeting of the rising flux tubes in the upper layers of the convection zone causes the scatter of the tilt angles (Longcope & Choudhuri 2002). This scatter in the tilt angles certainly introduces a randomness in the poloidal field generation process.

The poloidal field gets built up during the declining phase of the cycle and becomes concentrated near the poles during the minimum. The polar field at the solar minimum produced in a theoretical mean field dynamo model is some kind of ‘average’ polar field during a typical solar minimum. The observed polar field during a particular solar minimum may be stronger or weaker than this average field. The theoretical dynamo model has to be updated by feeding the information of the observed polar field in an appropriate way, in order to model particular cycles. We propose the following methodology for this in our recent paper (Choudhuri *et al.* 2007). We run the dynamo code from a minimum to the next minimum in the usual way. If we want to model a particular minimum with a known observed value of DM, then we multiply the poloidal field of the theoretical model stopped at the minimum by a constant factor everywhere above $0.8R_{\odot}$ such that its amplitude becomes proportional to the value of DM. Since some of the poloidal field at the bottom of the convection zone may have been produced in the still earlier cycles, we leave it unchanged by not doing any updating below $0.8R_{\odot}$. Only the poloidal field produced in the last cycle which is concentrated in the upper layers gets updated to make it correspond to the observed DM. After this updating, we run the code till the next minimum, when the code is again stopped and the same procedure is repeated by using the value of DM at that minimum. Our solutions are now no longer self-generated solutions from a theoretical model alone, but are solutions in which the random aspect of the dynamo process has been corrected by feeding the observational data of polar fields into the theoretical model.

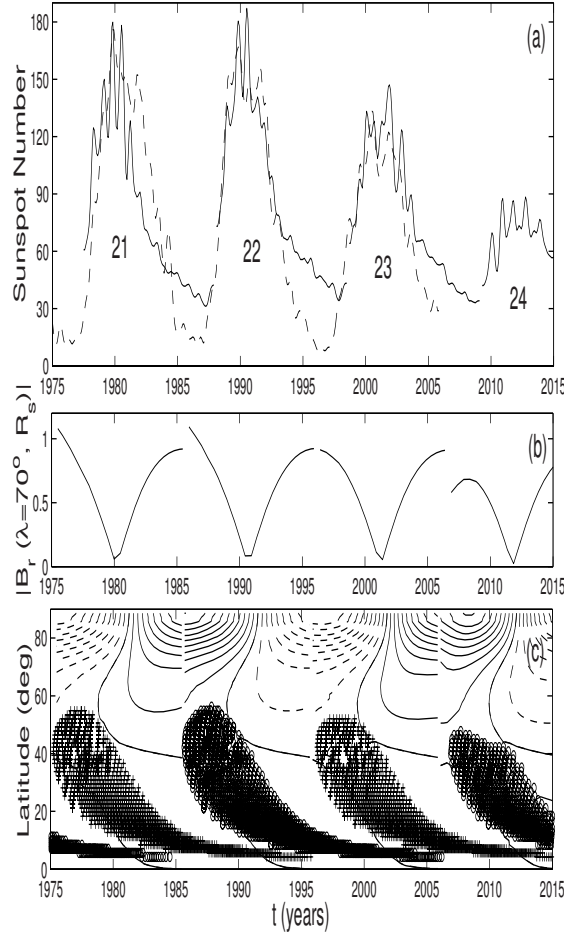


Figure 3. Theoretical results from our dynamo model for cycles 21–24. **(a)** The theoretical monthly smoothed sunspot numbers (solid line) superposed on the monthly smoothed sunspot numbers from observation (dashed line). **(b)** A plot of B_r at the surface at a latitude of 70° . **(c)** The theoretical butterfly diagram, with contours of B_r at the surface in the time-latitude plot. This figure is taken from Choudhuri *et al.* (2007).

Our calculations are based on a dynamo model presented in some of the earlier papers from our group (Nandy & Choudhuri 2002; Chatterjee *et al.* 2004). Since values of DM are available only from the minimum at the end of cycle 20 and afterwards, Choudhuri *et al.* (2007) could apply their methodology to model the last few cycles only. Figure 3 shows their results. Since the value of DM at the end of cycle 23 (i.e., at the present time) is available, calculations could be continued till the minimum at the end of cycle 24. The top panel in Fig. 3 shows theoretical sunspot numbers (solid line) superposed on the observed sunspot numbers (dashed line). We find that cycles 21–23 are modeled reasonably well by our solar dynamo model. Cycle 24 comes out much weaker than even the reasonably weak cycle 23, suggesting that it will probably be the weakest cycle in a century. The middle panel plots the radial magnetic field at a point on the surface. We see the discontinuities at the minima when the poloidal field is

updated in accordance with observational data. The bottom panel gives the theoretical butterfly diagram.

The result that cycle 24 will be a weak cycle seems like a very robust result from our dynamo model and does not change significantly on varying various parameters within reasonable limits. It may be noted that the strength of cycle 24 predicted by us would make it lie exactly on the straight line in Fig. 1 indicating the possible correlation.

3. Conclusion

Choudhuri (1992) postulated several years ago that stochastic fluctuations in the poloidal field generation mechanism may be the cause of irregularities in the solar cycle. We now identify the randomness in the Babcock–Leighton process as the source of these fluctuations. Because of this randomness, a strong cycle is not necessarily followed by a strong polar field at its end and vice versa. The lack of correlation seen in Fig. 2 can be taken as a confirmation of our hypothesis that the Babcock–Leighton process involves randomness. Although we have reliable polar field data only for the last 4 minima, the polar field at earlier times can be indirectly estimated from polar faculae data (Sheeley 1991) or positions of filaments (Makarov *et al.* 2001). These data also suggest that the strength of the polar field at a minimum determines the strength of the next cycle, but the strength of the cycle is not very well correlated with the polar field produced at its end.

Theoretical considerations alone do not allow us to build a model of actual solar cycles. We have to feed some relevant information from observational data into our theoretical model. Since the polar field produced at the end of a cycle cannot be predicted correctly from a mean field model (due to the randomness in the production process), we suggest that ‘correcting’ the polar field at the minima by using observational data is the best way of feeding relevant observational information into the theoretical model. In contrast, Dikpati & Gilman (2006) use the sunspot area data as the source of poloidal field in their model. In view of the randomness in the Babcock–Leighton process leading to a lack of correlation between the sunspots of a cycle and the polar field at its end (as seen in Fig. 2), we are highly doubtful whether the methodology of Dikpati & Gilman (2006) is correct. We shall have a verdict from the Sun-god himself within the next 4–5 years whether their prediction or our prediction comes closer to the truth.

The rising phase of a solar cycle is dominated by processes like the advection of the poloidal field by meridional circulation and its stretching by the differential rotation in the tachocline, which are ordered and deterministic processes. So we suggest that the rising phase of a cycle is reasonably deterministic, allowing us to predict the strength of a cycle a few years ahead of time by using the polar field data. On the other hand, the declining phase of a solar cycle is not predictable, since it is dominated by the Babcock–Leighton process which involves randomness. Hence, it may never be possible to predict the strength of a maximum more than 7–8 years ahead of time even when we have better dynamo models and more high-quality magnetic field data.

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What Helicity Can Tell Us about Solar Magnetic Fields

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Abstract. Concept of magnetic/current helicity was introduced to solar physics about 15 years ago. Earlier studies led to discovery of such fundamental properties as hemispheric helicity rule, and role of helicity in magnetic reconnection and solar eruptions. Later, the concept was successfully applied in studies of different solar processes from solar dynamo to flare and CME phenomena. Although no silver bullet, helicity has proven to be a very useful “tool” in answering many still-puzzling questions about origin and evolution of solar magnetic fields. I present an overview of some helicity studies and briefly analyze their findings.

Key words. Sun—magnetic fields—helicity.

*“If you are after good publicity,
you should not speak about current helicity.”*

From a poem written by Jan Stenflo

1. Knots and bolts

The concept of helicity has its origin in the knot theory. Since its development, it has been successfully used in different fields of mathematics, plasma physics, and more recently, in solar physics and astrophysics.

To demonstrate the usefulness of helicity concept in solving topological problems, let us consider a seemingly complex object – a tangled knot (Fig. 1a). Is this knot topologically similar to a simple O-ring? One approach would be to untangle the knot and demonstrate that indeed it can be transformed to a ring. This direct approach, however, might not be always practical especially if one has no physical access to an object. Alternatively, one can calculate helicity of the knot; if helicity $H = 0$, then the knot is topologically the same object as O-ring. For the purpose of this example, helicity can be calculated by the means of “helicity meter” (Fig. 1, right panel). In this exercise, a thread forming the knot should be followed in one (arbitrary selected) direction and every crossing should be counted as +1 or –1. When applying helicity meter, the top arrow should always be aligned with thread in the direction it is followed. Orientation of lower arrow will determine sign of helicity: positive +1 or negative –1. In the course of this exercise, one can find that the knot on Fig. 1 has equal number of positive and negative crossings, and hence, zero helicity, similar to ring.

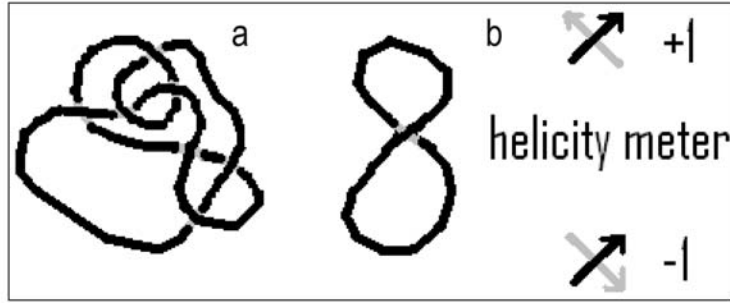


Figure 1. Knot (a) and figure-8, (b) as two examples of topological objects, which can be transformed to a simple O-ring. Right panel shows helicity meter that can be used to determine helicity of two objects shown to the left. Numbers next to helicity meter show corresponding helicity.

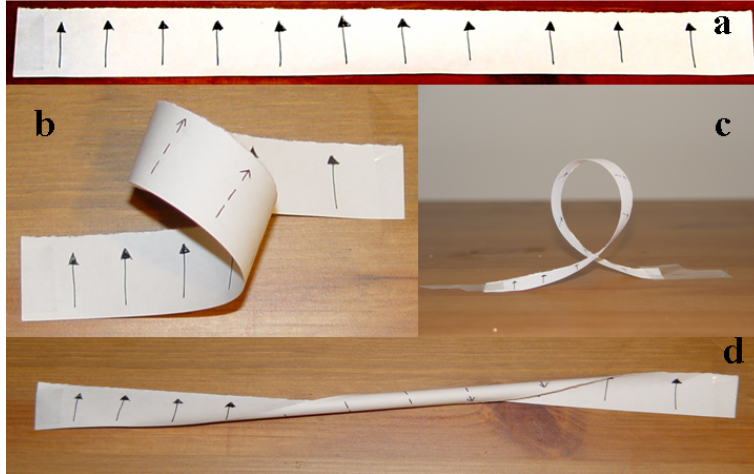


Figure 2. A ribbon (a) demonstration of two components of helicity: writhe (b and c) and twist (d).

However, counting crossings to determine helicity may be misleading. For example, folding O-ring on itself into figure-8, will not change helicity. However, helicity meter applied to Fig. 1(b) will yield, negative helicity, $H = -1$, contrary to the fact that figure-8 can easily be converted to O-ring. This is because helicity may have several components, while “helicity meter” can only be used to find one component.

Figure 2 demonstrates two components of helicity: twist, T_w and writhe, W_r . Wrapping a straight paper ribbon (Fig. 2a) around itself creates writhe (Fig. 2b and c), while pulling two ends of ribbon apart, transforms writhe to internal twist (Fig. 2d). Total helicity, H , of this ribbon, shown in Fig. 2(b–d) is $H = T_w + W_r$. Figure 2(c) corresponds to $T_w = 0$, $H = W_r = -1$ (using helicity meter). Figure 2(d) corresponds to $W_r = 0$, $H = T_w = -1$, and hence, left-hand twist corresponds to negative helicity.

Mathematically, helicity is defined as a dot product of a vector and its curl, integrated over a closed volume. Thus, for example, magnetic helicity, $H_m = \int \mathbf{A} \cdot \nabla \times \mathbf{A} dD$, where D is closed volume, and $\mathbf{A}$ is vector potential of magnetic induction $\mathbf{B} = \nabla \times \mathbf{A}$. Similarly, one can define current helicity, $H_c = \int \mathbf{B} \cdot \nabla \times \mathbf{B} dD$, and kinetic helicity, $H_k = \int \mathbf{V} \cdot \nabla \times \mathbf{V} dD$, where $\mathbf{V}$ is velocity of flows. Assuming that magnetic field is represented by a thin flux tube model with flux Φ , one can show that magnetic helicity, $H_m = (2\pi)^{-1} \Phi^2 (T_w + W_r)$ (Longcope and Klapper 1997), similar to our pictorial example on Fig. 2.

In a restrictive case of linear force-free field, $\nabla \times \mathbf{B} = \alpha \mathbf{B}$, $\alpha = \text{const}$, magnetic vector potential $\mathbf{A}$ can be expressed as $\mathbf{A} = \alpha^{-1} \mathbf{B} + \nabla \psi$, where ψ is an arbitrary scalar function. Then, magnetic helicity can be written as

$$H_m = \int \mathbf{A} \cdot \mathbf{B} = \int (\alpha^{-1} \mathbf{B} + \nabla \psi) \mathbf{B} dD = \int \alpha^{-1} B^2 dD + \int \nabla \psi \mathbf{B} dD. \quad (1)$$

For a closed volume, magnetic field does not cross the volume boundary, which results in additional condition $\mathbf{n} \cdot \mathbf{B} = 0$. Under this condition, the second integral goes to zero, and magnetic helicity can be expressed in terms of α -coefficient, and magnetic energy, E_m :

$$H_m = 2\mu\alpha^{-1} E_m, \quad (2)$$

where μ is magnetic permeability in vacuum, and $\alpha = \text{const}$.

Magnetic helicity has several properties making it an important parameter to study. Due to inverse cascading, magnetic helicity conserves better than magnetic energy, and it plays an important role in (or contains important information about) such processes as magnetic reconnection, dynamo, and stability of magnetic fields (see individual articles in Brown *et al.* 1999; Büchner and Pevtsov 2003).

2. Observations of helicity

To determine H_m requires the knowledge of magnetic field and its vector potential throughout the closed volume encompassing this magnetic field. On the Sun, however, magnetic field is only measured in a few levels (typically one or two) currently restricted to the photosphere and low chromosphere. The detailed distribution of subphotospheric magnetic field is unknown. Thus, to derive even limited information about helicity requires additional assumptions.

In the framework of force free field model, α coefficient has the same sign as magnetic helicity (see, equation 2). This sign correlation encouraged use of α as helicity proxy (e.g., Pevtsov *et al.* 1995). Using vector magnetic observations in a single level in solar atmosphere, one can determine α either by fitting linear force free field to the observed (transverse) field (so called, α_{best}), or by computing vertical, z , component of α for each pixel in magnetogram and averaging it ($\langle \alpha_z \rangle$). Burnette *et al.* (2004) found a reasonably good correlation between α_{best} and $\langle \alpha_z \rangle$. Alternatively, one can also compute z -component of current helicity density, $h_c = (\mathbf{B} \cdot \mathbf{J})_z = \mu B_z \cdot J_z$, where J is electric current density. Although current helicity is not a conserved quantity (unlike magnetic helicity), it has the same sign as H_m . Two different helicity proxies based on current helicity were used: a fractional imbalance of h_c (percentage of pixels of one sign of h_c in a given magnetogram) and area-averaged h_c ($\langle h_c \rangle$) (e.g., Abramenko *et al.* 1996; Bao and Zhang 1998).

It has also been shown that sign of helicity may be inferred from the morphological patterns observed in various solar phenomena. For example, spiral pattern of filaments forming sunspot (super-) penumbra may be interpreted as left-hand (LH) or right-hand (RH) twist. Orientation of barbs in chromospheric filaments indicates two distinct types of filaments: dextral and sinistral. Chirality of filaments is thought to be associated with the handedness of their magnetic fields. Shape of sigmoidal coronal loops can be directly related to $\alpha > 0$ (S-shape) or $\alpha < 0$ (N-shape) magnetic field. A relationship between these patterns and magnetic field helicity proxy α has been established at least in a framework of linear force free field model (see, individual articles in Brown *et al.* 1999).

One can also compute change of helicity relative to a reference, e.g., potential field, so called relative helicity (e.g., Chae 2001 and references therein).

3. Hemispheric helicity rule

All helicity proxies discussed reveal a common tendency that became known as the hemispheric helicity rule. Solar magnetic fields in the Northern Hemisphere have negative sign of helicity, while the fields in the Southern Hemisphere possess positive helicity. The hemispheric helicity rule was found in the magnetic fields at different spatial scales, from network magnetic fields, sunspots, large-scale fields, and magnetic clouds, from the photosphere, through the chromosphere, corona, and solar wind (Brown *et al.* 1999). Typically, about 70–80% of solar features (e.g., active regions) follow this hemispheric dependency. A pictorial summary of the hemispheric helicity rule can be found in Pevtsov and Balasubramaniam (2003).

Several mechanisms were considered as potential origin of this rule including solar differential rotation, direct action of the Coriolis force, solar dynamo, and turbulent convection in upper portion of the convection zone (Σ -effect). A summary of these mechanisms can be found in Longcope *et al.* (1999) and Longcope & Pevtsov (2003). Similar to Joy's law that describes active region tilt relative to equator, the hemispheric helicity rule shows significant scatter, suggesting that turbulence in the convection zone may have some effect. It now appears that interaction between magnetic field and turbulent convection in upper portion of the convection zone (so called Σ -effect) is the leading mechanism behind the hemispheric helicity rule. The mechanism was originally proposed by Longcope *et al.* (1998), and more recent studies found additional support to it (e.g., Nandy 2006).

Both historic and more recent studies suggest that the hemispheric helicity rule does not change from one solar cycle to the other. However, some researchers suggested that the rule might change its sign in some periods of solar cycle. Thus, for example, Bao *et al.* (2000) found reverse sign of the rule for h_c at the beginning of Cycle 23. On the other hand, α_{best} helicity proxy showed no change in the hemispheric helicity rule for the same period (Pevtsov *et al.* 2001). Hagino & Sakurai (2002) found indication that some periods of solar activity cycle disobey the helicity rule. Choudhuri *et al.* (2004) suggested a theoretical model that predicts deviations from the hemispheric helicity rule at the beginning of each solar cycle. On the other hand, Pevtsov *et al.* (2003a) pointed out lack of consistency between different magnetographs in respect to years when the hemispheric rule reverses its sign. They suggested that due to significant scatter in the data, years with low sunspot activity (e.g., beginning of solar cycle) may show a deviation from the rule simply because of insufficient statistical sample. One

should also mention a recent study by Zhang (2006) suggesting that helicity of umbral and penumbral fields may have opposite sign, and different solar cycle behavior. Thus, the question of cycle dependence of the hemispheric helicity rule remains open (see, Pevtsov *et al.* 2008).

4. Helicity transport

Whatever the mechanism that generates helicity on the Sun, it operates continuously. Thus, helicity should be removed from the Sun at approximately the same rate as it is created. Accumulation of helicity would make it difficult for solar dynamo to operate – a problem referred to as dynamo quenching (e.g., Brandenburg and Sandin 2004). The prime way for the Sun to shed helicity is via coronal mass ejections (CMEs), although some helicity may also be dissipated “on site” by reconnecting different flux systems that have opposite signs of helicity. These opposite helicity reconnection events are likely to release significant amounts of energy and play role in (some) solar flares (e.g., Kusano *et al.* 2004).

Exactly, how much helicity is removed by CMEs and is there sufficient helicity in solar magnetic fields to account for that amount?

Lepping *et al.* (1990) fitted 18 magnetic clouds (MCs) to a linear force-free field (LFFF) flux tube model (typical $\alpha \approx 10^{-10} \text{ m}^{-1}$, $B_0 = 2 \times 10^{-4}$ Gauss). In the framework of LFFF model, helicity of MC can be expressed as $H_{MC} = \alpha^{-1} L \Phi^2$, where L is total length of flux tube and Φ is its magnetic flux. The exact length of MC flux tube can be estimated to be between twice the distance to the Sun and an arch with its two footpoint anchored at the Sun. With these uncertainties, H_{MC} was found to be about $5 \times 10^{42} \text{ Mx}^2$. Demoulin *et al.* (2002) studied a long-term evolution of relative helicity in the active region NOAA AR7978. The region produced 26 CMEs over the period of one solar rotation, shedding $\approx 5.2 \times 10^{43} \text{ Mx}^2$ of magnetic helicity. However, magnetic field of the region persisted over five solar rotations continuing to produce CMEs and MCs. It was estimated that the total helicity ejected by MCs often exceeded coronal helicity of this source active region; and that differential rotation could not effectively replenish helicity lost to an MC.

Total magnetic helicity of active regions can be estimated using equation (2). For linear force-free field magnetic energy can be derived using Virial theorem on the basis of a single magnetogram:

$$E_m = \mu^{-1} \int (x B_x + y B_y) B_z dx dy, \quad (3)$$

where x, y are Cartesian coordinates, and $B_{x,y,z}$ denote two horizontal and the vertical component of magnetic field. This approach was tried on about 160 active regions observed in a chromospheric spectral line using National Solar Observatory at Kitt Peak full disk longitudinal magnetograph (Jones *et al.* 1992). Magnetic field of selected active regions was fitted by a set of LFFF models, and the resulting field lines were fitted to coronal images observed by EIT/SOHO in 195Å (Urdike & Pevtsov 2002). Resulting distribution of helicity of active regions is shown in Fig. 3. Mean helicity of active regions’ magnetic field is about $1.7 \times 10^{43} \text{ Mx}^2$, about three times larger than a typical helicity of a magnetic cloud.

Thus, it is clear that magnetic field of active regions has sufficient amount of helicity to support continuous helicity removal by CMEs. On the other hand, several studies

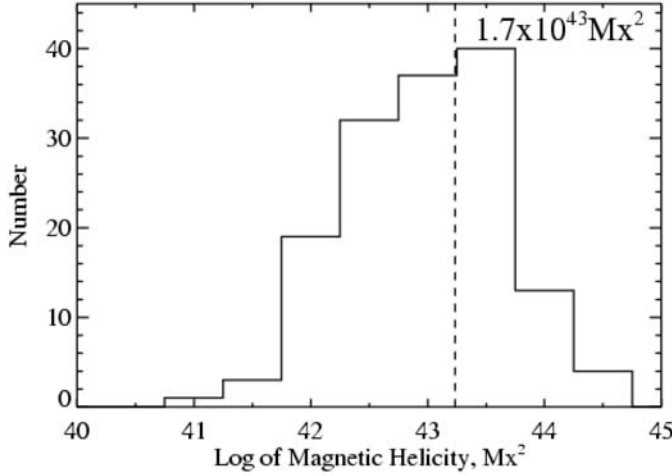


Figure 3. Distribution of magnetic helicity of active regions computed under the assumption of linear force-free field. Vertical dashed line shows mean of the distribution.

show that observed surface motions are insufficient to produce the required amount of helicity. The same studies suggest that helicity observed in the photosphere and corona originates below the surface. This raises that question, how helicity generated by subphotospheric processes is transported to the corona?

Longcope and Welsch (2000) considered evolution of twisted flux tubes as they rise through the photosphere to the corona. As flux tube crosses into the corona, it expands. This expansion disrupts torque balance between narrow and wide portions of the tube, and hence, some twist should be redistributed to the expanded portion of the tube. These authors predicted a particular behavior of twist depending on the rate the flux tube emerges. Pevtsov *et al.* (2003b) studied evolution of helicity proxy α in several emerging active regions, and found it to be consistent with Longcope and Welsch (2000) predictions. Fitting the model to the observations yields that subphotospheric twist will propagate from about 10 Mm to the corona in about one day. This suggests a possible mechanism of helicity transport from below the surface to the corona and (via CMEs) to the interplanetary space. When CME removes helicity from the corona above active region, helicity from the subphotospheric part of active region will replenish corona, and within a day the region will be able to produce another CME. Helicity will be transported by torsional waves, which may have very little observational signature in transverse displacements in the photosphere, which might explain the lack of significant horizontal motions that could be responsible/related to helicity buildup. Chae *et al.* (2003) considered a somewhat similar model of helicity pumping from narrow to wide portion of an expanded flux tube.

5. Some conclusions

Significant developments in the field of helicity makes it impossible to provide a comprehensive review of all aspects of helicity studies in solar physics in one short

article. From the topics that we had visited, however, one can make the following conclusions:

- Magnetic field on the Sun exhibits the hemispheric helicity rule. Deviations from the rule were observed to occur in some (early) years of solar cycle, but lack of consistency between the instruments calls for additional studies to confirm if the rule may vary with the solar activity cycle.
- Helicity of strong (active region) magnetic fields observed in the photosphere is probably generated in mid-upper convection zone. Pevtsov & Longcope (2001) observations also suggest that helicity of weak (network) field may originate in the convection zone; surface dynamo only recycles fields generated below the surface. These observational results need to be verified using high resolution magnetograms from Hinode.
- It is suggested that helicity is constantly transported to the corona via torsional motions. In this scenario, magnetic field below the surface may serve as an “untapped” reservoir of helicity, which can replenish corona within a day after a major CME eruption removes helicity from the Sun.

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The Evolution of Vector Magnetic Field Associated with Major Flares in NOAA AR10656

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Abstract. In this paper, we study the evolution of vector magnetic field of AR 10656 by using the observations of Huairou Solar Observing Station (HSOS, China) and Big Bear Solar Observatory (BBSO, USA). The magnetic flux emergence and cancellation, and thus, magnetic non-potential changes, are associated with the major flares in this active region. Compared with some other super-active regions, the evolution of magnetic morphologies and non-potentialities are relatively gradual, and thus the energy transportation and release are relatively slow. This gradual process may result in the recurrent flares of AR 10656.

Key words. Magnetic field—evolution—flares.

1. Introduction

Flares are believed to be powered by the magnetic energy stored in the nonpotential component of magnetic fields. This nonpotential component can be deduced from the difference between the observed vector magnetic field and the inferred potential field, which is calculated from the observed line-of-sight magnetic field (Rust *et al.* 1994). The causal relations between the non-potentiality and flare occurrence have been discussed by many authors. Wang *et al.* (1996) presented time sequences vector magnetograms obtained at Huairou Solar Observing Station (HSOS) and showed a great enhancement in the non-potential field several hours before an M-class flare. They also found that some decrease in the non-potential field during and after the flare. Vector magnetograms revealed line-of-sight electric currents before the flare. These currents disappeared after the flare. Based on the vector magnetograms observed at HSOS, Bao *et al.* (1999) showed that rapid and substantial changes of distribution of current helicity in an area or in its vicinity are most likely the triggers of flares. Also based on the vector magnetograms observed at HSOS, Deng *et al.* (2001) found that, for the “Bastille Day flare”, in the photosphere and before the flare, the distribution of angular shear underwent a dramatic change, and the vertical current system, current helicity and source field were disrupted and weakened. A recurrent solar flare activity

Table 1. Major flares occurred in NOAA AR10656 (from Space Environment Center).

Date	Begin	Max	End	Class	Date	Begin	Max	End	Class
12	0438	0505	0520	M1.2	15	0445	0506	0522	M1.2
13	0636	0729	0738	M1.2	15	0554	0600	0604	M1.2
13	1202	1209	1212	M1.1	15	1123	1132	1148	M2.6
13	1807	1812	1815	X1.0	15	1234	1241	1243	M9.4
13	2314	2343	2346	M3.0	15	1837	1845	1850	M1.2
13	2346	2350	2354	M2.8	16	0331	0347	0415	M1.1
14	0410	0414	0417	M2.4	16	2229	2244	2252	M1.1
14	0536	0544	0552	M7.4	17	1926	1937	1948	M2.4
14	0751	0756	0759	M2.3	17	2112	2121	2139	M1.8
14	0952	1007	1017	M3.2	17	2212	2228	2237	M1.3
14	1331	1343	1350	M5.6	18	1729	1740	1754	X1.8
14	1809	1818	1832	M1.3	19	0635	0701	0718	M3.0
14	2009	2016	2059	M1.3	19	1329	1351	1416	M2.1

including micro to major flares occurred in NOAA active region 10656 (S13 W36) between 08 and 18 August 2004 (Jain *et al.* 2007). An X1.0 flare exploded near the disk center on 13 August 2004, and X1.8 on 18. In addition, there were 24 M-class flares in this region during its disk passage (Table 1). The vector magnetic fields of this region were observed at HSOS and BBSO. Based on these vector magnetograms, and SOHO/MDI observations, we will study the daily evolution of vector magnetic fields and its association to the occurrence of flares in this active region.

2. Observations and data reductions

The vector magnetic field data used in this paper were obtained by the Solar Magnetic Field Telescope (SMFT) (Ai & Hu 1986; Deng *et al.* 1997) at HSOS and the digital magnetograph system (DMG) at BBSO. With the SMFT, photospheric vector magnetic field is measured with a Zeeman line FeI 5324.191 Å. The field of view and the spatial resolution of the data are about $5.23' \times 3.63'$ and $2-3''$ respectively. The noise of longitudinal and transverse fields are about 20G and 200G respectively. The CaI 6103 Å was selected in DMG (Cacciani *et al.* 1990), and in this system the spatial resolution and sensitivity can be markedly improved to $1''$ and 0.5 G (Spirock *et al.* 2001). The registration hardware is 512×512 , 12-bit CCD that allows 30 frames per second rate. White light and longitudinal magnetic images of MDI/SOHO (Scherrer *et al.* 1995) are also used in our analyses. The data reductions follow the methods introduced by Wang *et al.* (1996) and Deng *et al.* (2001). Here we only give the summaries in our data reductions. We use the linear force-free approximation to determine the 180° ambiguity. As to the inconsistency between the sensitivities of transverse and longitudinal magnetogram, we will not correct the projection effects. This is reasonable because this active region is quite close to solar disk center during our interested period.

3. The evolution of magnetic morphology

Figure 1 shows the evolution of sunspots and longitudinal magnetic fields in AR 10656, observed by MDI/SOHO. From 07 to 09, the region seems to be a simple bipolar

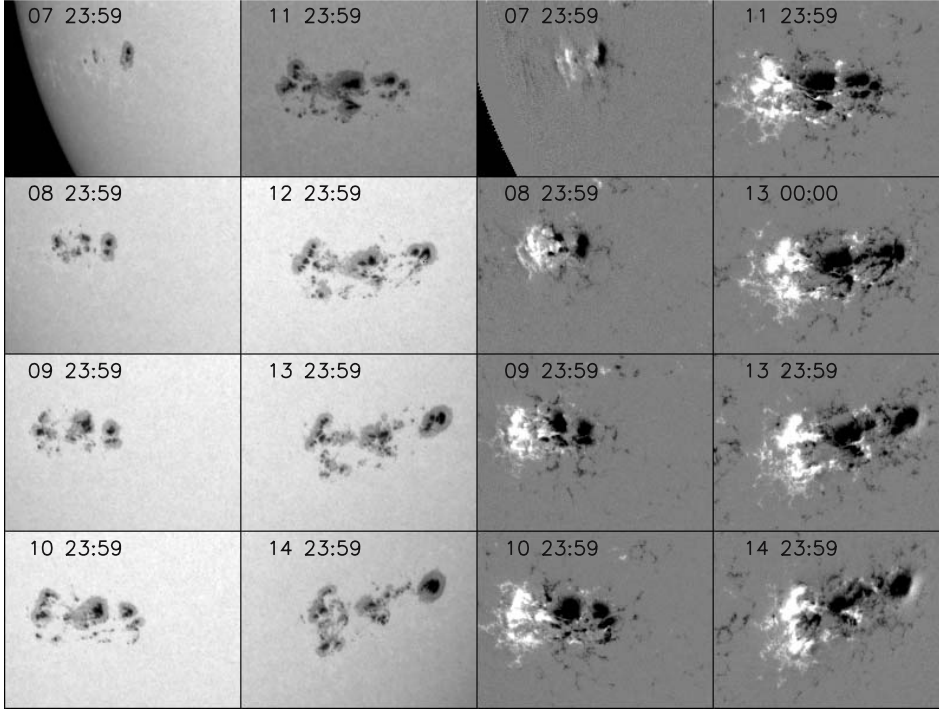


Figure 1. Daily evolution of sunspots (left panel) and longitudinal magnetic field (right panel) in NOAA AR 10656 from 07 to 14 August 2004.

sunspot. After that, new flux emerges between the leading and following sunspot. These emerging fluxes are found to increase in their intensity and area, move towards opposite polarity and finally cancel the flux. This region matches $\beta\tau\delta$ classification. Figure 2 shows the change of tilt angle, which is defined as the angle between magnetic axis of the active region and the local latitude line (Tian *et al.* 1999). There is about 20 degrees of counter clockwise rotation. As this active region is in south hemisphere, it matches Joy's Law. All these phenomena indicate new and rapid flux emergence, and as a result, frequent flare activities (as shown in Table 1).

Guo *et al.* (2006) suggested that the effective magnetic distance could quantify the magnetic configuration of active regions. After studying 43 active regions, they concluded that when the value of effective distance is more than 1, the active region is more complex and active, and when the value is less than 1, the active region is relatively simple and not active. The effective magnetic distance of AR 10656 is 1.03, which means that this region is quite active, but not superactive. Some major flares may happen, but the possibility of massive X-class flares is low.

4. Evolution of the nonpotentiality from 09 to 14 August 2004

Figure 3 gives the daily evolution of longitudinal current distribution. One can see that the current system is a bit stable and changed gradually and slowly during the six days. Some obvious changes occurred in part "A" and "B". Current system "A"

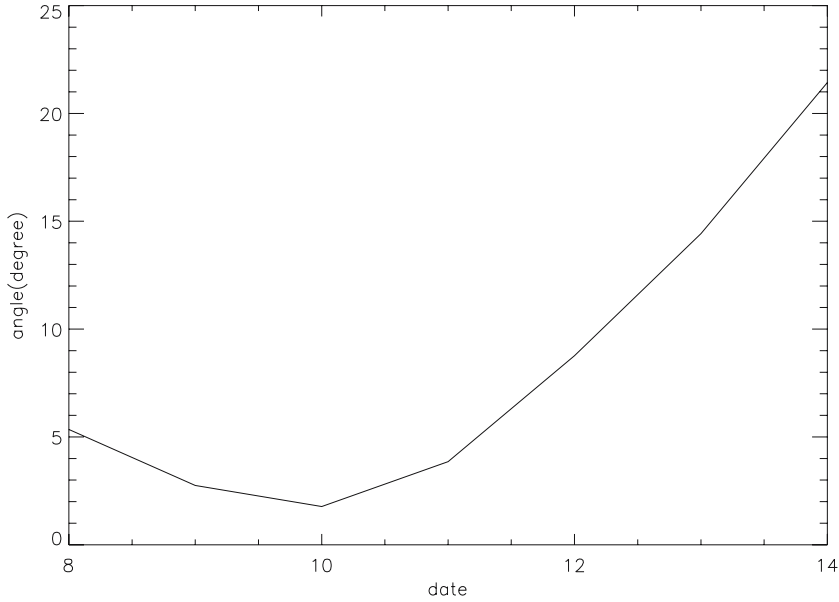


Figure 2. The change of tilt angle deduced from MDI/SOHO magnetogram.

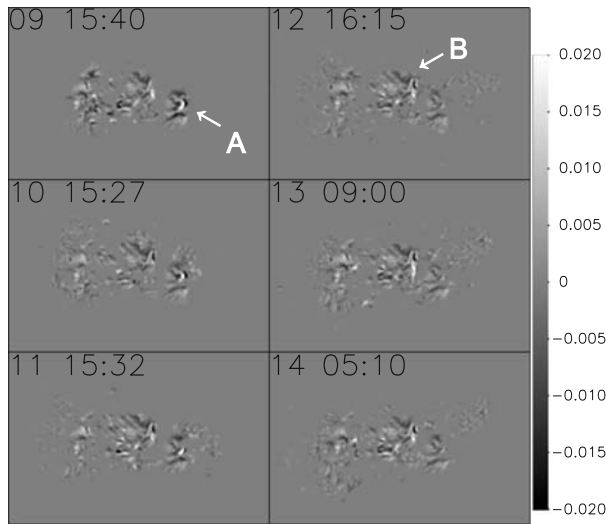


Figure 3. Daily evolution of longitudinal current from 09 to 14 August. The distribution of longitudinal current is scaled between $\pm 0.02 \text{ Am}^{-2}$. Data from 09 to 12 were collected with DMG/BBSO, the others by SMFT/HSOS.

gradually decreased from 09 to 11, and almost disrupted after 11. System “B” gradually increased, and reached maximum around August 13, and then tended to decrease. The current helicity exhibited a similar evolution pattern. There may be some relationships between these parts and the major flares. This will be discussed in another paper.

We also study the evolution of shear angle and free magnetic energy. Basically, the conclusion is similar with that from daily evolution view – no sudden change happened.

5. Conclusions and discussion

Based on the vector magnetograms observed at HSOS and BBSO, we have studied the evolution of magnetic nonpotentiality of AR 10656. This region has frequent new emerging flux and belongs to $\beta\tau\delta$ classification, and recurrent solar flare activity including micro to major flares occurred. Its effective magnetic distance is about 1.03, which implies that this region is active, but not super active, i.e., low possibility of X-class flares. Deng *et al.* (2001) found that before an X5.7 flare in AR 9077 there were sudden changes of vertical current, helicity, and free magnetic energy in one flare foot point (Box “B” in Deng’s reference). Liu *et al.* (2007) studied the nonpotential evolutions of 11 super-active regions, and found that for most super flare activities, obvious changes in nonpotentialities, such as current, helicity, shear, and free magnetic energy, can be found. Compared with these super solar flares, such as “Bastille” (Deng *et al.* 2001) and “Halloween” (Liu *et al.* 2007) events, AR 10656 has a relatively gradual daily evolution of magnetic nonpotentiality. We suggest that the magnetic topological structure is complex, but there is no heavy destruction of the magnetic morphology. Thus, the process of energy transportation and release from lower atmosphere to chromosphere and corona is relatively gradual. These may result in the recurrent flare activity of AR 10656, and low possibility of X-class flares.

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Inductive Magnetic Footpoint Tracking by Combining the Minimum Energy Fit with the Local Correlation Tracking and Doppler Velocity

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Abstract. Time-dependent magneto-hydrodynamic simulations of active region coronal magnetic field require the underlying photospheric magnetic footpoint velocities. The minimum energy fit (MEF) is a new velocity inversion technique to infer the photospheric magnetic footpoint velocities using a pair of vector magnetograms, introduced by Longcope (2004). The MEF selects the smallest overall flow from several consistent flows by minimizing an energy functional. The inferred horizontal and vertical flow fields by the MEF can be further constrained by incorporating the partial or imperfect velocity information obtained through independent means. This hybrid method is expected to give a velocity close to the true magnetic footpoint velocity. Here, we demonstrate that a combination of the MEF, the local correlation tracking (LCT) and Doppler velocity is capable of inferring the velocity close to the photospheric flow.

Key words. Velocity—magnetic field—induction equation.

1. Introduction

In the solar photosphere, it is now possible to measure all the components of the magnetic field with some accuracy. But at the same time only the line-of-sight component of the velocity is measurable at the photosphere. Hence, many techniques have been developed to infer the velocity of the structures in the solar atmosphere. Local Correlation Tracking (LCT; November & Simon 1988) is the most widely used technique to infer the velocity of features in the intensity images of the solar atmosphere. Application of this technique to follow the magnetic footpoints led to the problems that LCT does not follow the induction equation. To overcome this problem many other techniques have been developed. Those techniques use the induction equation and LCT velocity as a partial velocity information (Kusano *et al.* 2002; Welsch *et al.* 2004). Longcope (2004) developed a method, the minimum energy fit (MEF) which uses the induction equation and three components of the magnetic fields to infer three components of the velocity fields by minimizing the energy type of functional. In this paper, we show the test results of incorporating the partial velocity information such as the LCT and Doppler velocity into the MEF.

2. The MEF and its combination with the LCT and Doppler velocity

The measurement of the magnetic fields in the photosphere yields its three components. As the vertical derivative of the velocity and magnetic fields are not available yet, Longcope (2004) used the vertical component of the ideal induction equation to derive the velocity components. The method selects the one velocity from a set of multiple velocities by minimizing the energy-type of functional

$$W = \frac{1}{2} \int_M (|v_h - u_h|^2 + |v_z - u_z|^2) ds, \quad (1)$$

which is also consistent with the induction equation. v_h , v_z are the MEF generated velocity and u_h , u_z are the partial velocity information obtained through other means. There are four possibilities: (1) keep u_h and u_z zero, (2) keep u_z zero and incorporate u_h from the LCT, (3) keep u_h zero and incorporate u_z from Doppler velocity and (4) incorporate u_h and u_z from the LCT and Doppler velocity into the MEF. We call the first possibility as the MEF, the second possibility as a combination of MEF + LCT, the third as a combination of MEF + u_z and the fourth possibility as a combination MEF + LCT + u_z . In this paper, we test this new velocity inversion techniques on the simulated magnetogram and compare the results with the true velocities obtained from the simulation.

3. Test results

In order to test these velocity inversion techniques we apply it to the simulated vector magnetogram whose velocity field is well known. The synthetic magnetograms are derived from the time-dependent MHD simulation of flux emergence. More details about the physical parameters used in the simulation can be found in Welsch *et al.* (2007). A pair of vector magnetograms has been extracted from the simulated datasets with a time difference between the two magnetograms being $\delta T = 1256$ seconds and

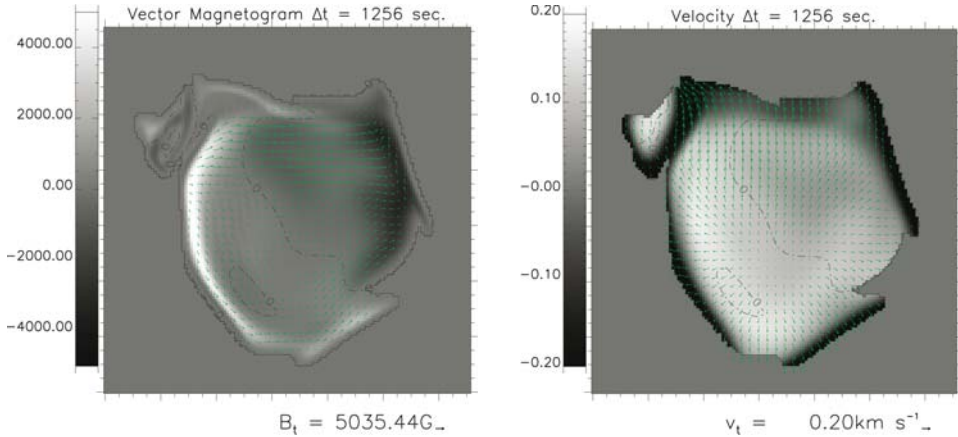


Figure 1. Left: A sample magnetogram. Right: A sample velocitygram extracted from the time series.

the pixel size 348 km. More detailed information about the extraction of the magnetograms and the time interval between them, etc. can be seen elsewhere (Welsch *et al.* 2007).

Figure 1 (left) shows the sample vector magnetogram. The magnetic field evolves with the velocity field shown in Fig. 1 (right). We applied our technique to those pixels whose magnetic field strength exceeds 170 G. To converge the MEF algorithm fast towards the smallest energy level we used a multi-grid technique. More details about our method can be found in Ravindra *et al.* (2007). We used a 15 pixel Gaussian apodizing window function for computing the LCT velocity and the obtained velocity is incorporated into the MEF. We extracted the vertical component of the velocity from the simulated data for that particular time interval and introduced into the MEF as u_z .

Figure 2 shows the inferred horizontal velocity vectors overlaid upon the vertical component of velocity (except in the case of LCT) obtained from the MEF, LCT,

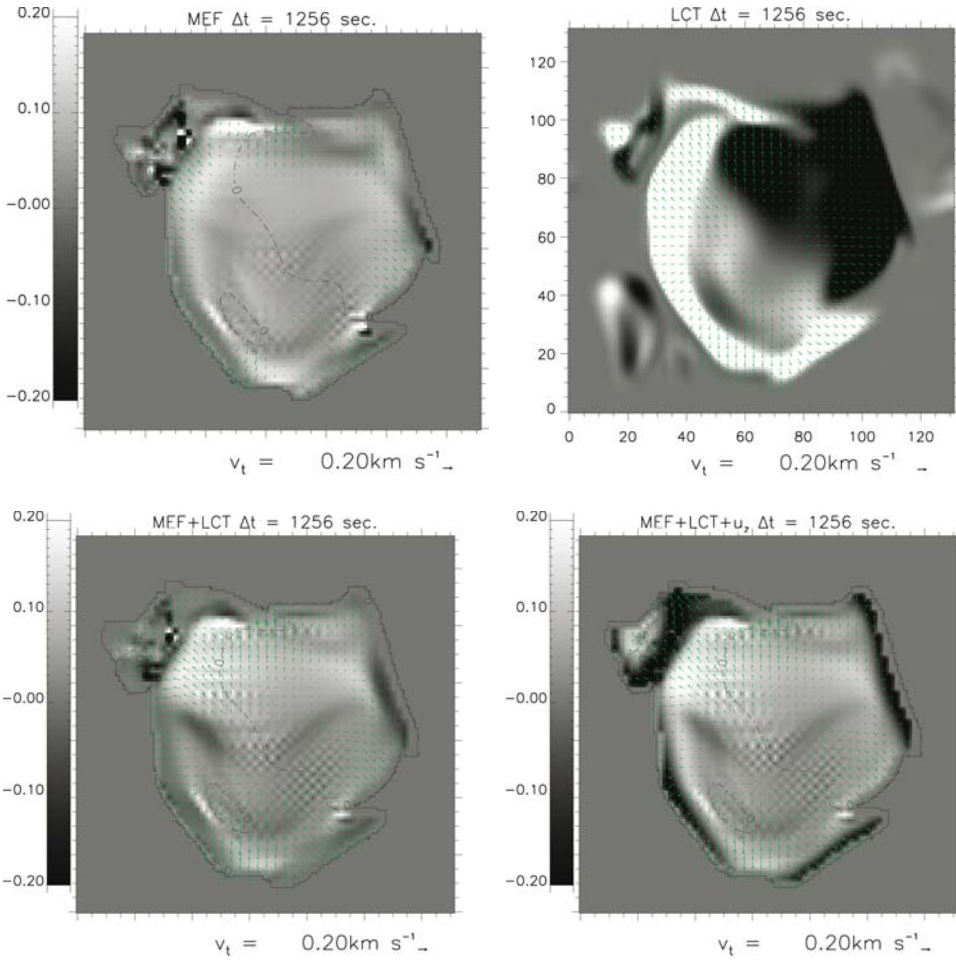


Figure 2. A horizontal velocity vector overlaid upon the vertical flow field obtained from MEF (top left), MEF + LCT (bottom left) and MEF + LCT + u_z (bottom right). Top right image is for LCT where the horizontal velocity vectors overlaid upon the magnetogram.

MEF + LCT + u_z and MEF + LCT (from top left in clockwise direction). In the case of LCT, the horizontal velocity vectors are overlaid upon the vertical component of the magnetic field as the LCT cannot measure the vertical component of the velocity. The direction of horizontal velocity vectors obtained from each of these techniques closely resembles the true flow field (Fig. 1 (right)). The vertical component of the velocity obtained from the MEF and MEF + LCT shows up-flow in most of the regions and little down-flows at the periphery. On the other hand, the vertical flows inferred from the MEF + LCT + u_z shows an up-flow in the center and down-flows in the periphery as seen in Fig. 1 (right). This is not a surprising result since the actual value of the vertical velocity is used as an input in the MEF + LCT + u_z algorithm.

Figure 3 shows the quantitative comparison of the magnitude of the inferred velocity using the MEF, LCT, MEF + LCT + u_z and MEF + LCT (from top left in the clockwise direction). These velocities are plotted against true velocity obtained by taking the average over a time difference of 1256 seconds which is the time interval between

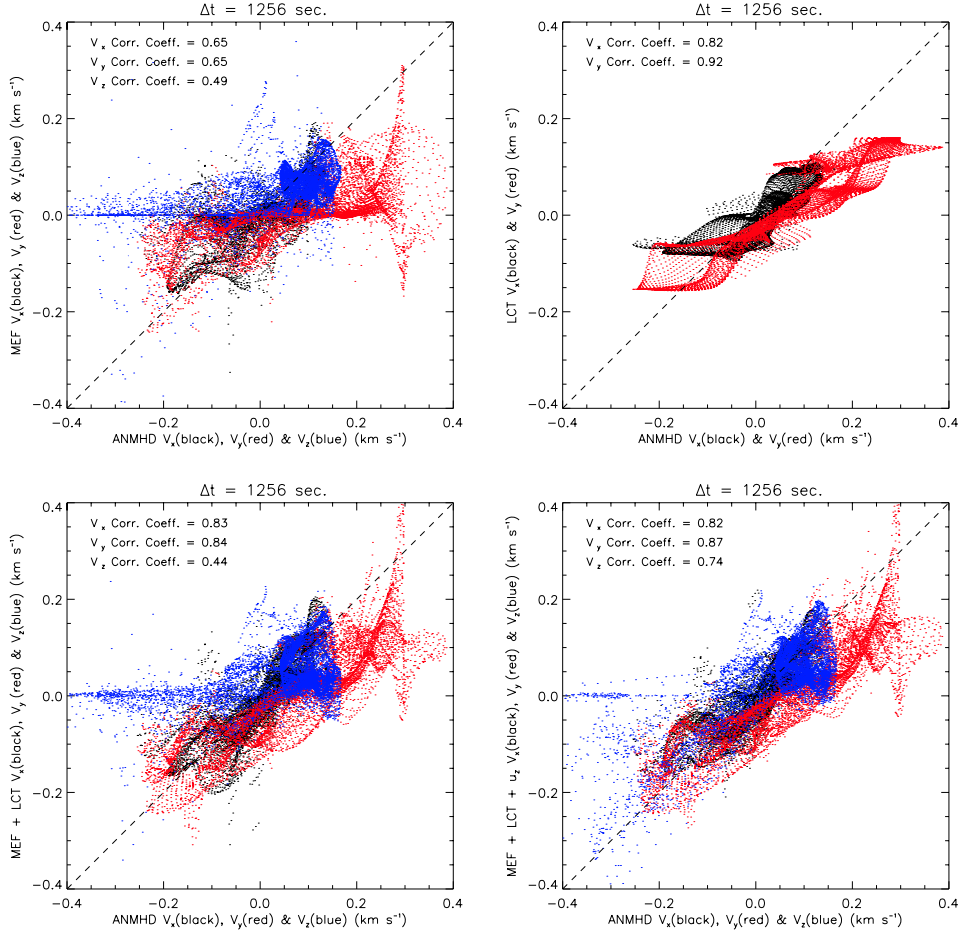


Figure 3. Scatter plot between the velocity vectors obtained MEF (top left), LCT (top right), MEF + LCT (bottom left) and MEF + LCT + u_z (bottom right).

the pair of vector magnetograms used in the test. A comparison of the velocity obtained from the MEF with the true velocity shows that there is a large scatter in their values. The magnitude and the correlation of horizontal velocity obtained from the LCT and combination of MEF + LCT improves. The magnitude and the correlation of the vertical component of velocity is best when we use the MEF + LCT + u_z algorithm. Also, the correlation between the horizontal velocity obtained from the MEF + LCT + u_z and the true flow field stands in between LCT and MEF + LCT.

4. Conclusions

The velocity inferred from the MEF in combination with LCT and Doppler velocity is consistent with the true flow fields. The direction and magnitude of the horizontal velocity obtained from each of these techniques closely resembles the true horizontal velocity with best correlation from LCT, MEF + LCT and MEF + LCT + u_z . The direction and magnitude of the vertical velocity is best inferred from the MEF + LCT + u_z technique. In the future, we would like to compare the other quantities like net vertical flow, Poynting flux and helicity flux estimated using the velocities obtained from each of these techniques with the quantities obtained using the true velocities. Finally, we would also like to apply this new velocity inversion technique to the real vector magnetogram data.

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Prediction of Peaks in Wolf Numbers in Cycle 24 according to Actual Numbers of Polar Faculae

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Abstract. Previous investigations by Makarov *et al.* have shown a relation between the peaks in the number of polar faculae and the peaks in the Wolf number. In cycles 20 and 22 the delay between peaks in polar faculae and Wolf number was 6.1 ± 0.1 year, north and south taken separately, as their peaks do not coincide. For the odd cycle 21, this shift was 5.6 years average. Polar faculae always *precede* the sunspots. The relevance of this for the dynamo mechanism is obvious. In cycle 23 the delay was 7.7 year (north) and 7.8 year (south). The approach of a deep minimum is probably responsible for this increased delay; thus for cycle 24 the delay between peaks of polar faculae and sunspots is expected to be at least 7.8 years and probably longer. The present polar faculae show 6 peaks above the smoothed average (north) and similarly 3 peaks south. The first peak for the sunspots will be at the earliest during the very end of 2007. As soon as one peak in the spots occurs the delay for cycle 24 can be estimated and the other peaks predicted.

Key words. Polar faculae—sunspots—dynamo—Wolf number.

1. Introduction

During a long time polar faculae (PF) were considered as a phenomenon *following* the sunspots (SP) like a tail end. Several publications by Makarov and coworkers (Makarov *et al.* 1989, 2008; Makarov & Makarova 1996) showed that the peaks in the number of PF preceded the peaks in the SP by about half a cycle. Makarov and Makarova (1996) found for cycles 19 to 21 an average correlation factor 0.84 between the PF and the SP following them (N and S separate), while the correlation between PF and the preceding SP was still 0.72 on the average. However, the meaning of these results is relatively restricted as both phenomena (PF and SP) wax and wane with the same period and roughly in a similar way. That the correlation between the PF and the following SP is higher than with the SP preceding them is due to the peaks and thus the peaks were considered separately, giving a clear correlation between PF and the SP following them and a bad correlation between PF and the SP preceding them.

On the other hand, the relation between the total number of PF and the total number of sunspots does not seem to be simple as the total number of PF increased during the

last 4 cycles by a factor larger than 3, while the number of SP decreased during the last cycle. We may interpret this situation at least partially as follows: the polar cap of the global magnetic field region increased during the last 12 cycles and this probably means that the region of PF is enlarged too, especially in the latitude band 35° – 55° where they are better observable. However, this can only be part of the explanation. It may be noted that the present tendency of the number of PF is again decreasing (PF cycle 24 will give about half of cycle 23), which suits the general tendency of approaching a deep minimum. Moreover, according to the Greenwich Observatory (1880–1954), the high-activity PF cycles of 1887 and 1944 were followed by very strong sunspot cycles, and the weakest PF cycles in 1879 and 1901 preceded the weakest sunspot cycles. A relation between the polar magnetic field and the sunspot number in the next sunspot cycle has been used for predictions of future sunspot numbers (Schatten *et al.* 1978; Schatten 1986; Makarov 2002; Duhau 2003; Svalgaard 2005). This was taken into consideration by some authors to predict the next cycle, e.g., Choudhury *et al.* (2007).

2. Cycles 20–23

Makarov *et al.* (2008) give figures and a table showing that for cycles 20 and 22 the peaks in the PF precede the peaks in the SP by 6.1 ± 0.1 year. For the odd cycle 21 it was 5.7 (N) and 5.4 (S) year. The correlation factor varied from 0.53 to 0.63. Peaks in the northern and southern hemisphere did not occur simultaneously but showed the same shift.

Figure 1(a and b) for cycle 23 confirms the correlation between the PF and the SP which follow the PF. However, now the shift is much longer: 7.7 (N) and 7.8 (S) year (correlation factor 0.61 (N) and 0.71 (S)). This contrasts even more with the short shift for cycle 21, i.e., the odd cycle clearly gave a shorter shift than the even cycles; in fact this suits the general feature that shorter PF cycles precede odd SP cycles (Makarov *et al.* 2003). The reason for the longer shift for cycle 23 is probably the approach of a deep minimum which according to Makarov *et al.* (2007) is expected to occur in cycle 26 ± 1 .

Another difference for cycle 23 is that the peaks north and south occur nearly simultaneously: about a month difference, while for cycles 20–22 this was several months. However, cycle 23 shows asymmetry between north and south in other aspects (Zharkova *et al.* 2004; Georgieva *et al.* 2007).

3. Cycle 24

Figure 2(a and b) represents the peaks of the PF, i.e., what remains after subtracting the smoothed average. The period is from 2001 till 2007, which is relevant for the SP in cycle 24. For the northern hemisphere we have 6 peaks (not counting the small ones), for the southern hemisphere we have 3 serious peaks. It is obvious that the N–S symmetry which occurred in cycle 23 has now disappeared. Taking the same shift of 7.8 years as for cycle 23 then we may expect peaks in the SP in 2008.0, 2008.9–2009.1, 2011.0, 2011.6, 2012.1 and 2014.2 (north) and 2010.5, 2011.5–2011.9, 2014.3–2014.5 (south) approximately. However, this shift is uncertain and it is probably longer. In that sense it seems necessary to wait for the first peak (or first 2 peaks) in SP to determine the shift which can then be used to predict the following peaks.

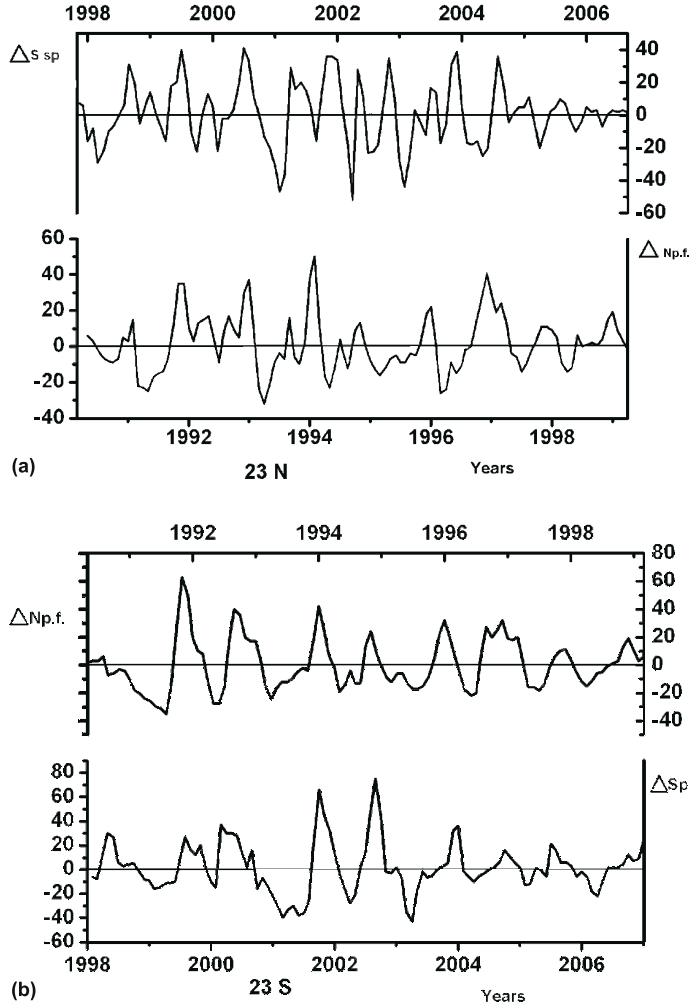


Figure 1. Cycle 23: the peaks in the number of polar faculae preceding the peaks in sunspots by 7.7 years (a) north and 7.8 years (b) south.

4. Conclusion

The fact that the PF precede the SP (at least the peaks do) has relevance for the Babcock-Leighton theory and for the theories of Choudhuri, of Dikpati and of Callebaut & Khater (2006). The latter theory has the advantage that it is based on an exact solution of the MHD equation and that PF and SP are generated naturally by the same mechanism, but its explanation (based on centrifugal forces) for the long delay of about half a cycle between PF and SP is now withdrawn. Instead, it is suggested that the meridional motion, on its return near the tachocline, releases first the PF and half a cycle later (and closer to the equator) the SP, which have been further amplified meanwhile and thus are stronger than the PF. This would solve the problem of the PF preceding the SP for all the theories mentioned above. The fact that the total number of PF per cycle is not

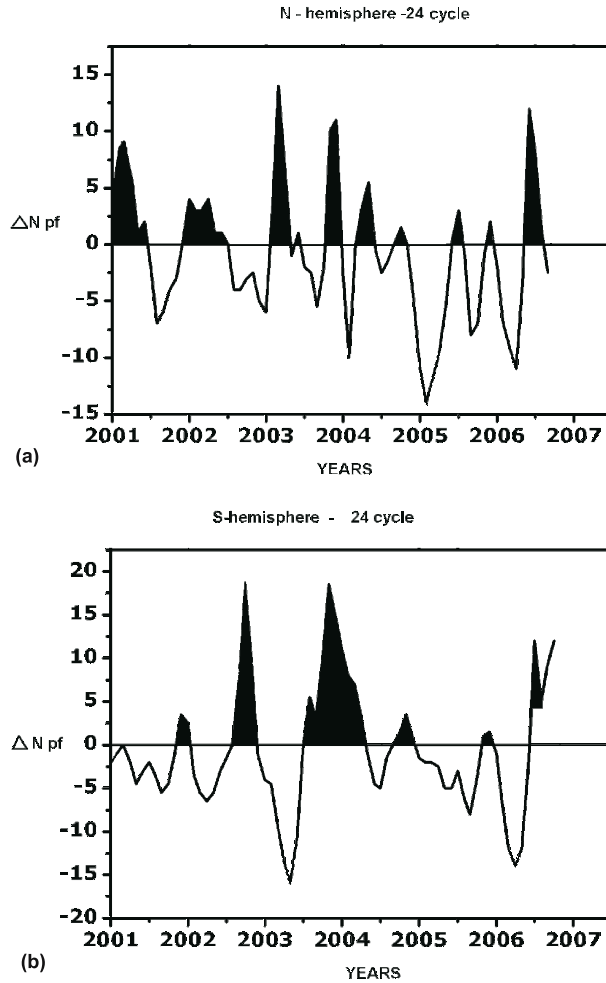


Figure 2. The peaks in the number of polar faculae (i.e., after subtracting the smoothed number of polar faculae): (a) north and (b) south. Period from 2001–2007, i.e., bearing on cycle 24 for the sunspots.

connected in a direct way with the total number of SP indicates that other effects have to be taken into account too.

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Helioseismology and the Solar Cycle: Past, Present and Future

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Abstract. A major goal of helioseismology is to understand the mechanism of the solar cycle. In this paper, some results of helioseismic observations relevant to the cycle are briefly reviewed, the current state-of-the-art is discussed, and near-term future directions are sketched out. Topics covered include the internal rotation rate; activity-related parameter variations; the tachocline; far-side imaging; the torsional oscillation; and meridional flows.

Key words. Helioseismology—solar activity cycle—tachocline—far-side imaging.

1. Introduction

Helioseismology provides information about the solar interior using precise measurements of the properties of the acoustic oscillations trapped within the Sun. One of the goals of helioseismology is to understand the mechanism of the solar activity cycle. Three solar cycles have now passed since it was observationally confirmed that the solar five-minute oscillations are trapped below the photosphere (Deubner 1975) and that the modes can be used to probe the solar interior. In addition, the modern era of helioseismology, which dates from 1995 with the deployment of the Global Oscillation Network Group (GONG) instruments and the launch of the Solar and Heliospheric Observatory (SOHO) space mission, has recently completed coverage of cycle 23. It is thus timely to review the subject.

In this paper, some of the helioseismic results pertaining to the solar cycle will be briefly discussed. The current state of our knowledge will be covered, and some speculations about the future direction of the field will be presented. Due to space limitations, this is not a complete review by any means, thus the author apologizes in advance to those authors whose work is not mentioned.

2. The past

Three topics are reviewed in this section: the discovery of activity-dependent frequency shifts; the determination of the internal solar rotation and the discovery of the tachocline; and the birth of local helioseismology.

Using integrated-light irradiance measurements from the Solar Maximum Mission, Woodard and Noyes (1985) found that the low-degree oscillation frequencies were

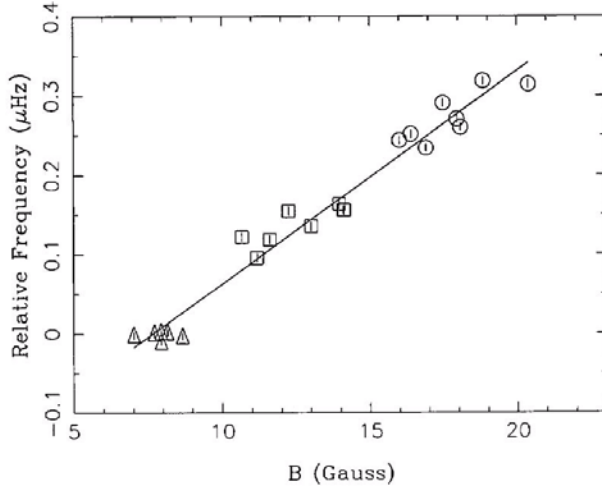


Figure 1. The linear relationship between frequency shift $\Delta\nu$ and surface magnetic field strength B as determined from Big Bear data in 1986 (triangles), 1988 (squares) and 1989 (circles). Adapted from Woodard *et al.* (1991).

0.4 μHz higher at solar maximum compared to solar minimum. This was the first evidence that oscillations varied with solar cycle. Subsequent work by Libbrecht and Woodard (1990) demonstrated that the frequency shifts were present for intermediate-degree modes, and had a strong frequency dependence. These observations were interpreted as evidence for latitude-dependent changes in solar structure associated with the activity cycle.

Woodard *et al.* (1991) studied the variations of the frequency on time scales of 22.8 days, one year and two years. They found that the frequency shift, $\Delta\nu$, is related to the surface mean magnetic field, B , by a linear relationship with a slope that is independent of the time interval between measurements (Fig. 1).

Since the observations show that $\Delta\nu$ depends strongly on frequency ν and only weakly on degree l , the source of the shifts must be close to the upper reflection point of the modes located immediately below the photosphere. It will be shown later that the spatial distribution of the shifts are highly correlated with the surface magnetic field, and that the slope $\partial\Delta\nu/\partial B$ depends on the time interval considered.

Prior to 1988, dynamo theories of the solar cycle (e.g., Gilman 1986) rested on the assumptions that the solar rotation rate decreased with depth from the surface, and that isorotation contours were parallel to the solar rotation axis (Fig. 2A). These assumptions were reasonable given the Taylor–Proudman theorem and the surface differential rotation. In 1988, the first 1.5-dimensional inversion of rotational splittings of global oscillation modes was published (Libbrecht 1988). The picture of the internal rotation was very different from that needed for the contemporary dynamo theories. The isorotation contours were roughly parallel to radii, not to the rotation axis (Fig. 2B).

The inversion also revealed the tachocline, the shear layer where differential rotation transitions to solid-body rotation (Fig. 2C). Located at base of convection zone, the tachocline is $\approx 0.05 R_{\odot}$ in thickness. The surface rotation rate matches the sub-tachocline rate at a latitude of about 40° . This is near the latitude of sunspot

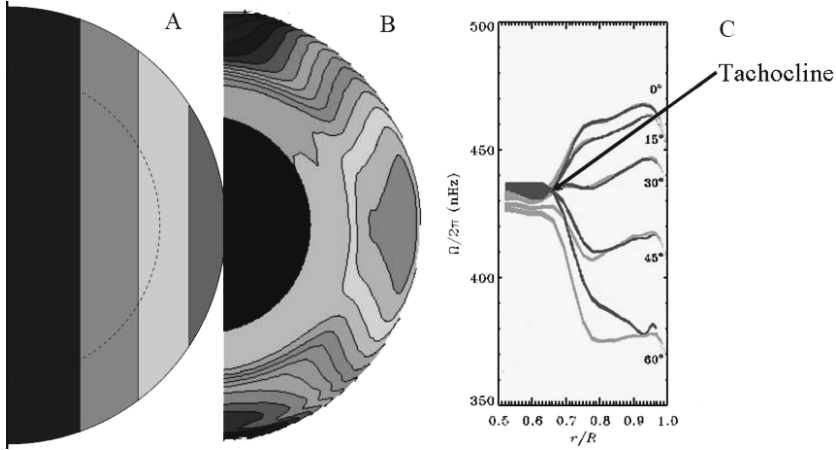


Figure 2. Panel (A): The isorotation contours assumed by pre-1988 dynamo theories. Panel (B): The isorotation contours revealed by helioseismology. Panel (C): Depth profiles of the rotational velocity at five latitudes. Dark grey lines are from MDI data, light grey lines are obtained from GONG data. The width of the lines shows the uncertainty in the measurements. The tachocline is indicated, where differential rotation in the convection zone transitions to solid-body rotation in the radiative envelope.

emergence early in the activity cycle. The existence of the tachocline was totally unexpected. It is currently thought that the dynamo operates in the tachocline, since a shear layer is an essential component of the dynamo mechanism.

The results discussed so far have been from the realm of global helioseismology, which decomposes the normal standing modes on the basis of spherical harmonics. This produces results that are averaged over the entire Sun, have no longitudinal information, and are symmetric in latitude across the equator. Since the Sun is not axisymmetric, it is of considerable interest to have helioseismic methods that can provide localized information. Such methods are known as local helioseismology. The common aspect of the techniques is the consideration of traveling waves rather than the normal modes. The analysis of the waves in localized areas can be used to obtain spatially resolved results. A comprehensive review of local helioseismology can be found in Gizon & Birch (2005).

Historically, the first local helioseismic analysis was the study of the absorption of p modes by sunspots. Braun *et al.* (1987) decomposed a circular area containing a sunspot into Hankel functions. They found that the amplitude of ingoing waves is higher than that of outgoing waves. This was the first indication that p modes can be used to sense localized conditions.

The ring-diagram analysis was developed in 1988 (Hill 1988, Morrow 1988). This analysis is based on a plane-wave representation and decomposes the waves in a localized area on the basis of sinusoids, valid for high-degree modes. In the resulting 3-D power spectrum at constant ν the ridges appear as rings, hence the name of the method. The position of the rings is a weighted average over depth of the horizontal velocity field below the analysis area. Thus, a 3-D map of the horizontal velocity field can be constructed.

Acoustic holography was developed by Lindsey and Braun (1990, 1997). In this method, a Green's function is computed that describes how an impulse originating from

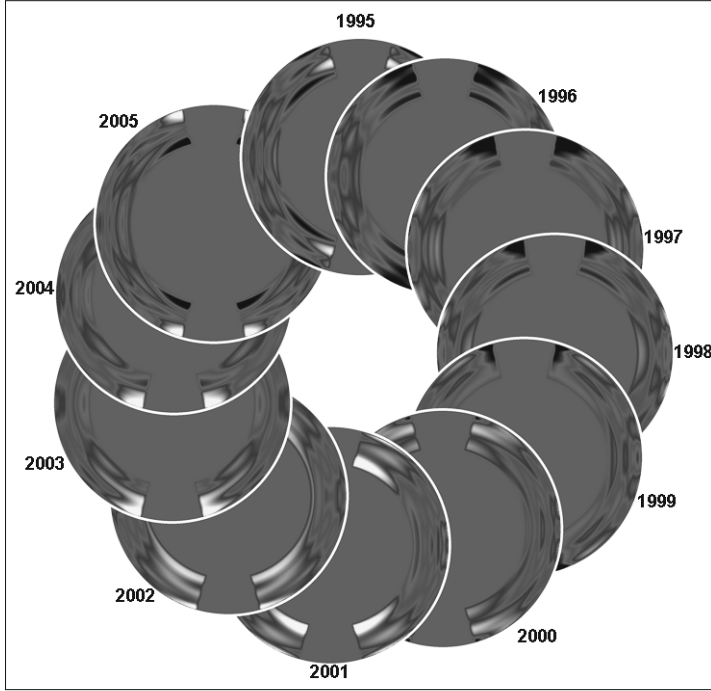


Figure 3. One solar cycle of convection zone velocity evolution as inferred from global helioseismic analysis of GONG data. Each circle is a cross-section of the solar interior for one-year intervals. Dark grey tones represent slower than average velocities, light shades are faster than average.

an arbitrary point appears at the surface. Using the Green's function, the wave field is mathematically time reversed to come to a focus at the coordinates of the original impulse. The focus can be located on the far side of the Sun as will be discussed later.

Time-distance analysis was initially developed by Duvall *et al.* (1993) and is very similar to methods used in terrestrial seismology. Acoustic waves emitted at the surface travel downward into the solar interior, are refracted upwards and reappear some time later and some distance away from the source. The time and distance is influenced by the conditions of the plasma that the wave propagates through. Time-distance diagrams are constructed by computing cross-correlation at many times T between a point and many distances D .

3. The present

We now turn to the present status of helioseismic analyses of the solar cycle. Space limitations restrict this section to only a few topics: convection zone changes; the tachocline oscillation; the torsional oscillation; surface distribution of mode parameter variations; far-side imaging; and meridional flows.

Recently, GONG and SOHO/MDI completed one solar cycle of imaging helioseismic observations covering cycle 23. Using the global modes, Howe *et al.* (2005)

have mapped the flows in the convection zone over the entire cycle. Figure 3 shows the evolution as a set of yearly averages of the velocity residuals. Several features are apparent. There are substantial fluctuations in the rotation rate in the tachocline. The torsional oscillation is clearly visible. Finally the polar regions are seen to rotate relatively slowly at solar minimum and accelerate as the cycle progresses. As the next minimum approaches, the high-latitude deep convection zone region decelerates. The deceleration region rises to the surface to complete the cycle.

Analysis of the residual rotation rate around the depth of the tachocline by Howe *et al.* (2000a) showed an apparent oscillation with a period of 1.3 yr visible in both GONG and MDI data. However, Antia and Basu (2000) were unable to confirm the result with an independent analysis. A more recent analysis by Howe *et al.* (2007) shows that the character of the oscillation may change during the cycle. Figure 4 shows the tachocline rotation rate as a function of time, and power spectra of the rotation rate for various epochs. From these spectra, it appears that there may be a coherent 1.3-yr oscillation in the rising phase that evolves into a more chaotic pattern in the declining phase.

The torsional oscillation was discovered by Howard and Labonte (1980) and appears as zonal flow on the surface with a velocity about 5 m/s greater than the local rotation rate. Active regions tend to emerge at the equatorward boundary of the flow. Work by Howe *et al.* (2000b, 2006) and Vorontsov *et al.* (2002) has demonstrated that the torsional oscillation pattern extends much deeper than the surface. The pattern is visible down to at least 40% of the depth of the convection zone, and there are indications that it may well exist throughout the entire CZ.

As discussed earlier, the frequency shifts associated with the solar cycle are mainly a surface effect. This is further demonstrated when the global frequency shifts are reprojected onto the solar surface. As shown in Fig. 5(A), the location of the shift regions is highly correlated with the activity bands (Howe *et al.* 2002). Studies of frequency shifts using ring diagrams also show large shifts, as much as 40 μ Hz, associated with specific active regions (Hindman *et al.* 2000). Finally, the lifetime and energy of the modes also vary with the activity level and are tightly correlated with the surface activity (Komm *et al.* 2002) as seen in Figs. 5(B) and 5(C).

Acoustic holography can produce maps of the phase shift arising from sound speed perturbations caused by surface activity on the far side of the Sun (Lindsey and Braun 2000). These far-side images can thus provide a prediction of the existence of large active regions as much as two weeks before solar rotation carries them to the front side where flares can influence terrestrial technology. Both GONG and MDI are currently providing continual near-real-time far-side maps on the web (González Hernández *et al.* 2006a). Figure 6 shows an example of a sequence of far-side maps showing AR10808 which subsequently produced a large X-class flare when it appeared on the east limb.

As discussed earlier, dynamo theories had to be revised after helioseismology revealed the true nature of the internal rotation rate. Modern flux-transport kinematic dynamo theories require a deep meridional return flow to transport flux from the poles back towards lower latitudes (e.g., Dikpati and Gilman 2006). So far, helioseismology has only been able to approximately measure the magnitude and depth of the return flow, the best current values from time-distance are 3 m/s at a depth of $0.8 R_{\odot}$ (Giles *et al.* 1997).

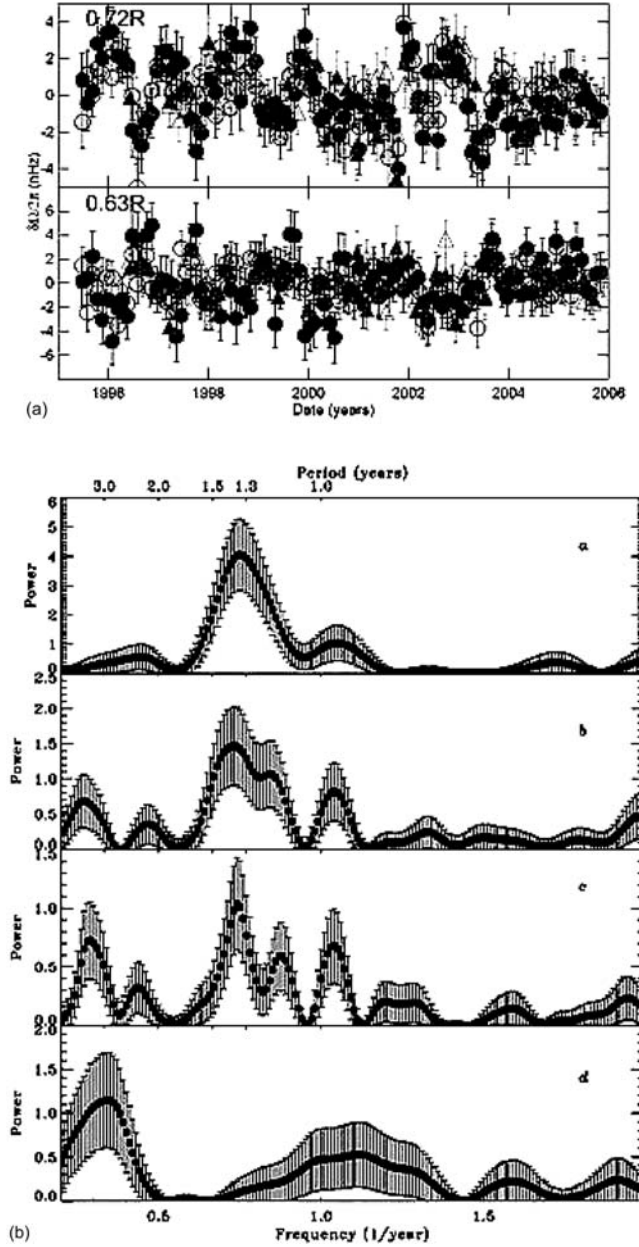


Figure 4. Variations in the rotation rate at the tachocline. (a) Residuals at the equator at $0.72 R_\odot$ (top) and $0.63 R_\odot$ (bottom) for RLS (filled) and OLA (open) inversions of GONG (triangles) and MDI (circles) data. (b) Power spectra of the rotation residual at depth $0.72 R_\odot$ and latitude 0° from GONG RLS inversion. Four time periods are shown: (a) 1995–2000, (b) 1995–2003, (c) 1995–2005, and (d) 2000–2005.

In addition, some numerical models of solar convection exhibit a two-cell meridional flow with a counter-cell at high latitudes (e.g., Brun and Toomre 2002). Analysis of

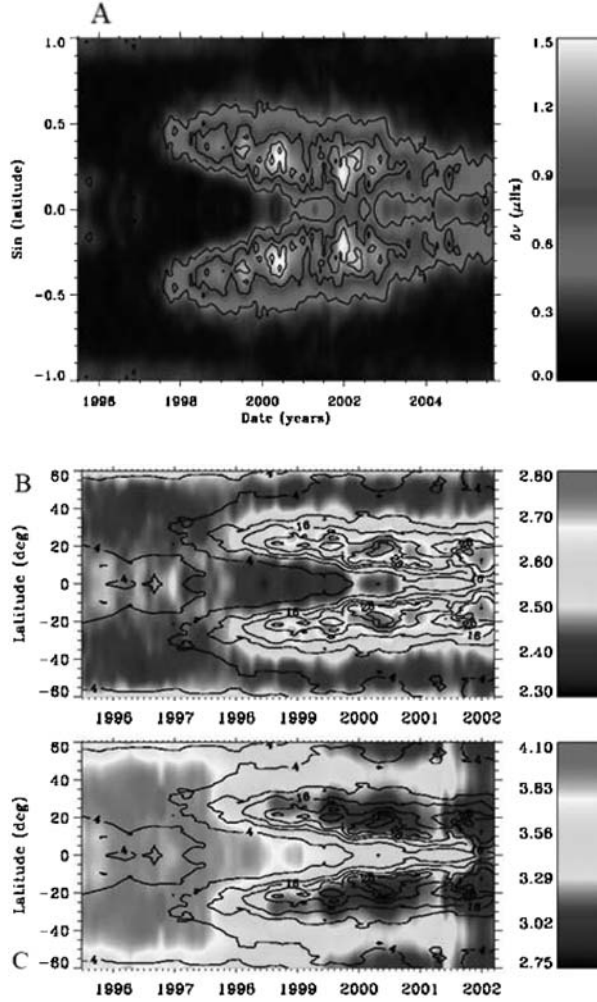


Figure 5. Global mode parameter shifts reprojected onto the surface show a very high correlation with surface magnetic activity. Panel (A): Frequency shifts. Panel (B): Mode width. Panel (C): Mode energy.

MDI data showed such a counter-cell near the north solar pole (Haber *et al.* 2002), but this is now thought to be a consequence of the annual variation of the solar B_0 angle (González Hernández *et al.* 2006b).

4. The future

What will be learned during cycle 24 from helioseismology? While predicting the future is always at best an approximation, there are some foreseeable developments. Since cycle 24 is the second half of a Hale cycle which began with cycle 23, there is considerable interest in continuing many of the current analyses to compare the internal solar dynamics between the two cycles. Will any difference be seen that can be related to the global polarity reversal?

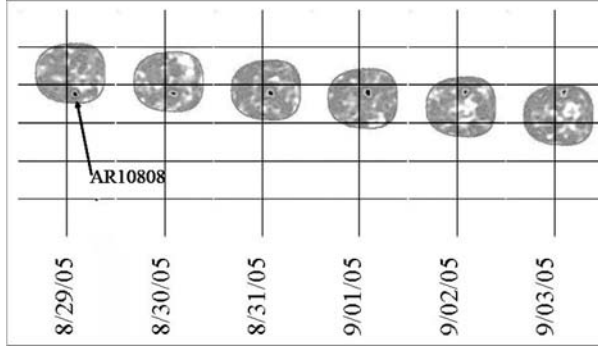


Figure 6. A sequence of far-side maps from 29 August 2005 to 3 September 2005 showing the signature of AR10808 as it rotated across the far side of the Sun. This active region produced a large X-class flare as it appeared on the east limb.

Of particular interest will be the behavior of the tachocline oscillation, if it exists. Will the 1.3-yr oscillation return with the ascending cycle? Another major question for cycle 24 is the existence of spatial structure in the tachocline. Can we see any signal of a structure in the tachocline that then evolves into an emerging active region? This will be a challenging observation for helioseismology – global analyses cannot provide spatial resolution, and the depth of the tachocline makes it difficult to observe with local methods.

Estimates of the characteristics of the deep meridional return flow are important to test modern dynamo theories. While some information will be gained by comparing the kinematic flux-transport dynamo predictions for the amplitude and timing of cycle 24 with what actually happens, it would be better to have helioseismic estimates. Current theories imply that the return flow has a magnitude of 1–2 m/s at the base of the convection zone, just beyond the capabilities of current helioseismic methods.

While the far-side images are already useful for predicting the appearance of active regions, they will be even more valuable if they can be calibrated in terms of active region properties. Recent work (González Hernández *et al.* 2007) has produced the first such calibration between the far-side signal, magnetic field strength, and area of active regions, but this initial work needs to be tested and improved.

The variation of the mode parameters over the solar cycle was the first indication that the oscillations can provide information about surface activity. It may thus be surprising if there is anything new to learn in this area, but recent work using very short nine-day time series has revealed a more complex relationship between the mode frequencies and the surface field. Tripathy *et al.* (2007) have shown that the slope of the linear relationship between $\Delta\nu$ and various activity indices varies with time and the phase of the solar cycle, as seen in Fig. 7. This suggests that there are two activity components that contribute to the variability of the slopes. The major contribution comes from the overall weak magnetic field, and hence during the ascending and descending phase of the cycle there is a strong correlation with activity. But during the maximum phase of the cycle, when the field is dominated by the strong field, due to the increase in the number and strength of the active regions, the slopes do not follow the activity measurements. Additional support for this scenario is provided by

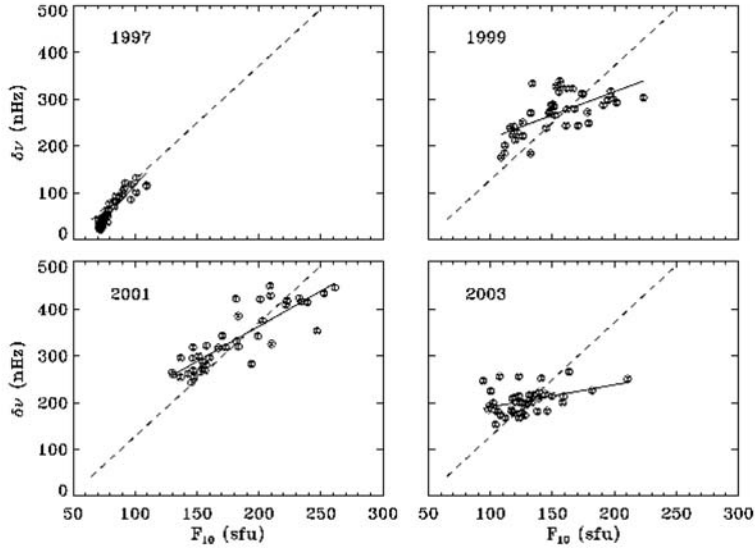


Figure 7. Mean frequency shifts plotted as a function of activity for four separate years. The solid line in each panel is the linear fit for that year. The dashed line is the linear fit for all datasets covering cycle 23.

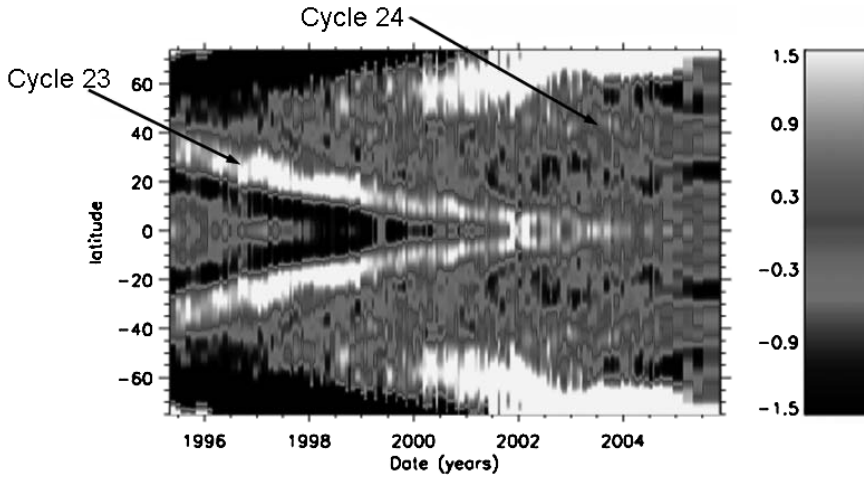


Figure 8. The torsional oscillation at a depth of 1000 km as function of latitude and time. The grey-scale bar indicates the rotation rate residual in nHz; $1 \text{ nHz} \approx 5 \text{ m/s}$. The flow associated with cycle 24 appeared in mid 2002, and seems to be substantially weaker than the flow for cycle 23. Does this predict a low amplitude for cycle 24?

similar results when the data are separated on the basis of latitude. Recent work by R. Howe (private communication) suggests that $\partial\nu/\partial B$ is different above and below 60° in latitude. Since no active regions are present at high latitudes, this may indicate that the interaction between the modes and the surface magnetic field is sensitive to the type of field.

Finally, every paper about the solar cycle should contain an audacious prediction about the next cycle. Figure 8 shows the temporal evolution of the torsional oscillation pattern at a depth of 1000 km below the photosphere. The flows associated with cycles 23 and 24 are indicated by arrows. The cycle 24 torsional oscillation appeared in mid 2002, and so far it seems that it is substantially weaker than the flow for cycle 23. Does this mean that cycle 24 will have a lower amplitude than cycle 23? Only time will tell.

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Seismic Study of Magnetic Field in the Solar Interior

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Abstract. Magnetic field in the solar interior contributes to the even order splitting coefficients, but it is not possible to separate the effect of magnetic field from those due to other deviations from spherical symmetry. Results obtained using GONG and MDI data are discussed. Limits on possible magnetic field in the solar core and in the tachocline region are obtained. There is some signal from possible magnetic field in the convection zone, but evidence of possible temporal variation in the solar interior is only marginal.

Key words. Sun: oscillations, interior, magnetic field.

1. Introduction

During the last decade detailed observations of solar oscillations have provided a unique tool to study the structure and dynamics of the solar interior. The frequencies of solar oscillations are characterised by the three quantum numbers, the radial order n , the degree l and the azimuthal order m . If the Sun was spherically symmetric the frequencies would be independent of m . Departures from spherical symmetry due to rotation, magnetic field or other sources lift the degeneracy and give rise to frequency splittings in modes of same n, l . The mean frequency of the multiplet is determined by the spherically symmetric structure of the Sun while the frequency splittings are determined by rotation, magnetic field and any other aspherical perturbations to the solar structure. To the first order rotation affects only the odd splitting coefficients, i.e., odd function of m and hence its effect can be separated from those of magnetic field and other asphericities, which contribute only to even splitting coefficients. The mean frequencies have been successfully used to study the structure of the solar interior (Gough *et al.* 1996) while the odd splitting coefficients have been used to study the rotation rate as a function of radius and latitude in the solar interior (Thompson *et al.* 1996; Schou *et al.* 1998). The inferred rotation rate can be used to compute the second order contribution from rotation to the even order splitting coefficients. These can then be subtracted from the observed coefficients to get the effect of magnetic field and other contributions to asphericity. Unfortunately, it is not possible to distinguish between the contributions of a magnetic field and aspherical perturbation to the solar structure.

The even order splitting coefficients are fairly small, and it is difficult to detect any signature of possible magnetic field in the solar interior. Dziembowski & Goode (1989) claimed to find evidence for a mega Gauss field near the base of the convection

zone, using data from the Big Bear Solar Observatory. Improved data from the Global Oscillation Network Group (GONG) project (Hill *et al.* 1996) and the Michelson Doppler Imager (MDI) instrument on board SOHO (Rhodes *et al.* 1997) have not confirmed these results (Antia *et al.* 2000). Instead of looking for signature of the magnetic field, we can assume that the even order splitting coefficients arise from asphericity in the solar structure. In this case, it is possible to apply an inversion technique to determine the latitudinal distribution of the sound speed and density (or other independent structure variables) in the solar interior (Antia *et al.* 2001). Of course, there is no way to distinguish whether the actual signal is due to magnetic field or asphericity. However, the advantage of this approach is that it is possible to identify the location of perturbation which is responsible for the even splitting coefficients.

The GONG and MDI instruments have been observing the Sun more or less continuously for the last 11 years and hence it is also possible to study possible temporal variation in the internal magnetic field or asphericity. It is well known that the frequencies of solar oscillations vary with time and this variation is correlated with the solar activity indices (Elsworth *et al.* 1990; Libbrecht & Woodard 1990; Bhatnagar *et al.* 1999). Similarly, the even splitting coefficients are also known to vary with time and their variation is correlated to the corresponding component of observed magnetic flux at the solar surface (Libbrecht & Woodard 1990; Woodard & Libbrecht 1993; Howe *et al.* 1999; Antia *et al.* 2001). However, most of these temporal variations are found to arise from perturbations near the solar surface (Basu & Antia 2000; Antia *et al.* 2001). There is only marginal evidence for any significant temporal variation in the solar interior. It is generally believed that the solar dynamo operates in the tachocline region and it would be interesting to study the possible presence of magnetic field and any associated temporal variations during the solar cycle.

2. Effect of magnetic field on solar oscillation frequencies

Since the forces due to rotation and magnetic field are much smaller than those due to gravity and pressure gradient, we can treat rotation and magnetic field as small perturbations over the spherically symmetric structure. This gives rise to splitting of frequencies with same n, l but different m . It is convenient to define the splitting coefficients by

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

where, $\nu_{n,l}$ is the mean frequency of the multiplet, $a_j^{n,l}$ are the splitting coefficients and $P_j^{n,l}(m)$ are a set of orthogonal polynomials of degree j in m (Ritzwoller & Lavelly 1991; Schou *et al.* 1994). The first order effect of rotation due to the Coriolis force affect only the odd order splitting coefficients, $a_1, a_3, a_5, \dots$ and these have been used to infer the rotation rate in the solar interior. The even splitting coefficients arise from second order effects of rotation, through the centrifugal force and magnetic field or other departures from spherical symmetry. Since forces due to rotation or magnetic field in the Sun are smaller by about 5 orders of magnitude as compared to gravitational forces, it is possible to apply a perturbative treatment to calculate their contribution to frequency splittings (Gough & Thompson 1990). The rotation rate in the solar interior can be inferred from the odd splitting coefficients and this can be used to estimate

its contribution to the even splitting coefficients. This estimated contribution can be subtracted from the observed coefficients to get the residuals which are due to magnetic field or other aspherical perturbations.

Following Gough & Thompson (1990) we consider a toroidal magnetic field of the form,

$$\mathbf{B} = \left[0, 0, a(r) \frac{dP_k(\cos \theta)}{d\theta} \right], \quad (2)$$

with the axis of symmetry coinciding with the rotation axis. Here, $P_j(x)$ is the Legendre polynomial of degree j and

$$a(r) = \begin{cases} \sqrt{8\pi p_0 \beta_0} \left(1 - \left(\frac{r-r_0}{d}\right)^2\right) & \text{if } |r - r_0| \leq d \\ 0 & \text{otherwise,} \end{cases} \quad (3)$$

where p_0 is the gas pressure, β_0 is the ratio of magnetic to gas pressure, which is assumed to be constant, r_0 and d are constants which determine the mean position and half-thickness of the layer where the field is concentrated. Using this magnetic field we calculate all second order contributions to the frequency splittings, which include the direct effect due to the Lorentz force and the contribution due to distortion from spherical symmetry introduced in the equilibrium state.

The splitting coefficients calculated from a given magnetic field configuration can be compared with observed splittings. We use the datasets from both GONG and MDI for this purpose. These datasets provide the mean frequency and the splitting coefficients for each n, l multiplet. We use 110 sets of GONG data each covering 108 days, starting from May 7, 1995 to April 14, 2006. The MDI data consist of 49 sets each covering 72 days, starting from May 1, 1996 to May 20, 2006. There was a break in MDI data between July 1998 and January 1999, when contact with SOHO was lost.

Since solar dynamo is believed to operate in the tachocline region we first consider magnetic field concentrated in the tachocline to calculate the splitting coefficients. These show a distinct signature in modes with lower turning point near the tachocline region. No such signature was found in observed splitting coefficients (Antia *et al.* 2000) and an upper limit of 300 kG in the tachocline region was estimated from these data. Even if the upper limit is reduced with accumulation of more data, it does not provide any significant constraint on theories of solar dynamo. The dynamo theories generally assume that the magnetic field is concentrated in flux tubes with a small filling factor, while in this work we have assumed an average field filling the whole tachocline region. If the dynamo field in flux tubes is averaged over the relevant region, its magnitude will decrease significantly, depending on the filling factor.

2.1 Magnetic field in the solar core

We can calculate the splitting coefficients due to a magnetic field in the solar core, but only a few modes penetrate to this region and the splitting coefficients of the low degree modes that penetrate to the core have large errors. Hence, it is difficult to make out any signal of possible magnetic field in the splitting coefficients. However, most of the solar

mass is in the core and any magnetic field in the core will cause significant distortion which would be visible even at the solar surface, unless it is compensated by a suitably large distortion in the outer layers. Using the formalism of Gough & Thompson (1990) it is possible to calculate the distortion at the solar surface due to large scale magnetic field in the interior. If we assume that the magnetic field is given by equation (2) with $\beta_0 = 10^{-4}$, $k = 2$ and $d = 0.1R_\odot$, we can calculate the resulting oblateness at the solar surface for different values of r_0 . When the field is located in the core, the resulting distortion at the surface is comparable to β_0 and as the field is moved outwards the distortion decreases. The estimated distortion is 2×10^{-4} , 6×10^{-5} , 10^{-5} , 6×10^{-6} for $r_0 = 0.2R_\odot$, $0.4R_\odot$, $0.6R_\odot$, $0.8R_\odot$, respectively (Antia 2002).

This distortion can be compared with observed distortion at the solar surface of $-(5.4 \pm 0.5) \times 10^{-6}$ (Kuhn *et al.* 1998). The seismically inferred rotation profile yields a distortion of -5.8×10^{-6} (Antia *et al.* 2000). Thus the unaccounted distortion at the surface is less than 10^{-6} . This can be compared with the expected distortion due to magnetic field to set an upper limit on the magnetic field in the core. The magnetic contribution will of course, depend on the form of the magnetic field and the region where it is located. Allowing for some uncertainties in that we can put a conservative upper limit of $\beta_0 = 10^{-5}$ in the solar core. This translates to a magnetic field of 7 MG at the centre, 3 MG at $r = 0.2R_\odot$ and 0.8 MG at $r = 0.4R_\odot$.

3. Asphericity in the solar structure

Apart from magnetic field and rotation, other departures from spherical symmetry in solar structure can also contribute to the even order splitting coefficients. Unfortunately, it is not possible to distinguish between these. Of course, the presence of magnetic field would yield some asphericity in solar structure, but that contribution is included in the magnetic field effects considered in the previous section. In this section we neglect the magnetic field and assume that even order splitting coefficients are due to asphericity in solar structure. Since the solar convection zone is not in hydrostatic equilibrium, such asphericities could arise due to anisotropy in turbulence. The advantage of this approach is that we can apply the inversion techniques to localise the perturbation (Antia *et al.* 2001). The splitting coefficients are sensitive only to the north-south symmetric component of asphericity and hence only this component can be determined. The differences in the sound speed, c , and the density, ρ , with respect to a spherically symmetric solar model can be expressed as (Antia *et al.* 2001)

$$\begin{aligned} \frac{la_{2k}^{nl}}{v_{nl}} = & \frac{Q_{lk}(4k+1)}{2} \int_0^R dr \int_0^\pi \sin \theta d\theta \left(\mathcal{K}_{c^2, \rho}^{nl} \frac{\delta c^2}{c^2} + \mathcal{K}_{\rho, c^2}^{nl} \frac{\delta \rho}{\rho} \right) P_{2k}(\cos \theta) \\ & + Q_{lk} \frac{F_k(v_{nl})}{E_{nl}}, \end{aligned} \quad (4)$$

where E_{nl} is the mode inertia (Christensen-Dalsgaard 2002) and Q_{lk} is a geometric factor as defined by Antia *et al.* (2001). Here $F_k(v)$ are the surface terms which accounts for uncertainties in the treatment of surface layers. This would also include contributions from outer layers which cannot be resolved by the mode set that is used for inversion.

Equation (4) can be used for inversion to determine the sound speed and density as a function of radial distance and latitude. Further, using data collected at different times,

it is also possible to study the temporal variations in these quantities. It is well-known that the mean frequencies as well as the even order splitting coefficients vary with time and it would be interesting to study if these imply any variation in solar structure with solar cycle. To bring out the temporal variations in the sound speed it is convenient to subtract the temporal average at each latitude and radius to obtain the residual which is the temporally varying component.

Figure 1 shows the results for aspherical component of $\delta c^2/c^2$ obtained from GONG data at $r = 0.96R_\odot$ as a function of time and latitude. There is no clear pattern in the residuals, thus suggesting that there is no significant temporal variation in sound speed. Similar results are obtained at other depths and it appears that most of the temporal variations in the even splitting coefficients is due to the surface term. Figure 2 shows

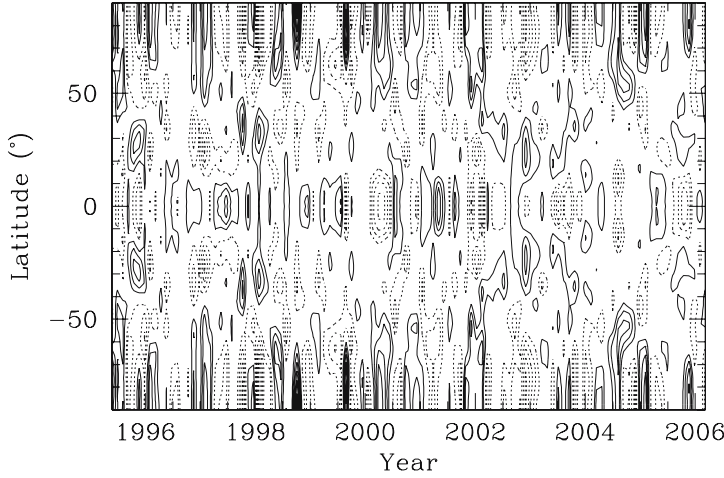


Figure 1. The contours of constant residual $\delta c^2/c^2$ at $r = 0.96R_\odot$ as a function of time and latitude for GONG data. Solid contours represent positive (greater than average) values, while dotted ones show negative values. The contour spacing is 2×10^{-5} .

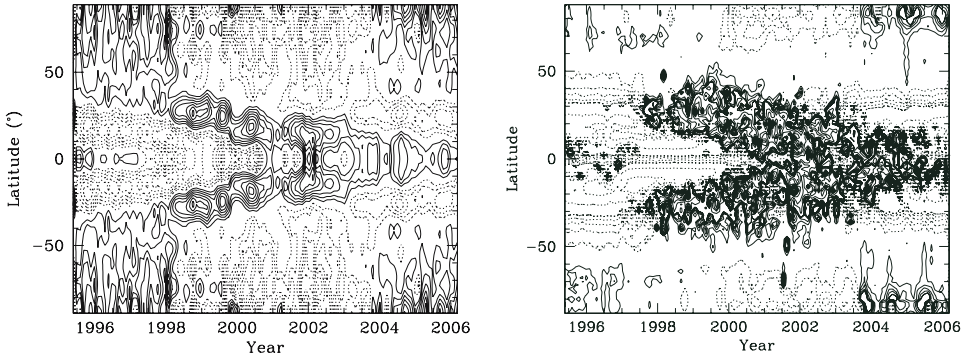


Figure 2. The surface term in asphericity inversion from GONG data (left panel) is compared with surface magnetic flux (right panel). Solid contours represent positive (greater than average) values, while dotted ones show negative (less than average) values.

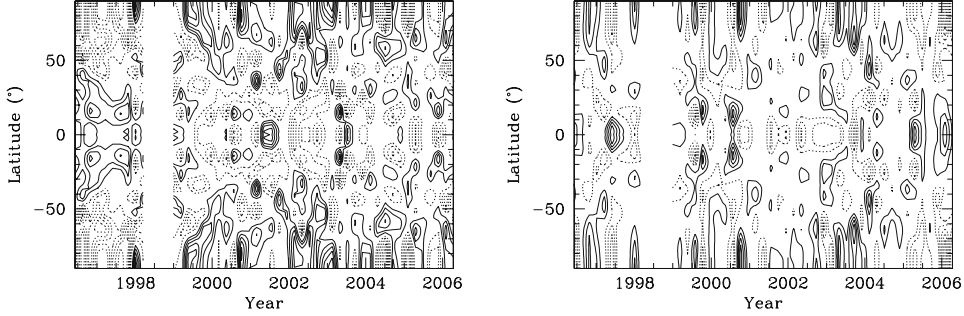


Figure 3. The contours of constant residual $\delta c^2/c^2$ at $r = 0.96R_\odot$ as a function of time and latitude for MDI data. The left panel shows the results using all modes, while the right panel shows the results using only modes with $l < 110$. Format is the same as that in Fig. 1.

similar variation in the surface term compared with that in the observed magnetic flux at the solar surface. It is clear that the surface term is well correlated with the observed magnetic flux at the solar surface. Thus it is quite likely that observed asphericity near the solar surface is actually due to magnetic field.

The left panel in Fig. 3 shows $\delta c^2/c^2$ obtained using MDI data. These results clearly show some temporal variations, but a closer look suggests that most of the variation has occurred during the data gap. Before the gap $\delta c^2/c^2$ is positive at low latitude and negative at high latitudes, while after the gap the situation is reversed. There are two possibilities, either the Sun had some interesting transition exactly when MDI was not operational or the difference is due to some instrumental variation during the recovery of the satellite. Since the GONG data do not show this variation, it is likely to be due to instrumental effect. This is further confirmed by the right panel, which shows the same results obtained from MDI data using only the modes with $l < 110$, which doesn't show much variation. Only high l modes appear to be affected by this artifact. These systematic errors in MDI data are likely to affect other results too (e.g., Antia 2003).

Since there is no significant temporal variation in asphericity, we can take temporal average over all datasets to improve accuracy and the results using GONG data are shown in the left panel of Fig. 4. MDI data give similar results if we restrict to modes with $l < 110$. The typical error in these results is about 10^{-5} , which is the contour spacing. The errors increase with depth and latitude. The figure shows a broad peak around $r = 0.9R_\odot$ and latitude of 60° . The maximum asphericity in c^2 is about 6×10^{-5} . This is well inside the convection zone and its origin is not clear. If this peak is due to magnetic field, then to an order of magnitude we can expect $\delta c^2/c^2 \approx v_A^2/c^2$, where v_A is the Alfvén speed. This would translate to a magnetic field of about 55 kG at $r = 0.9R_\odot$. It is not clear how such a field would survive inside the convection zone. This asphericity or magnetic field doesn't appear to have significant variation with solar activity.

Although there is no clear temporal variation in asphericity, in order to improve the statistics we can take temporal averages during the high and low activity phases and look at the difference between these. The results are shown in the right panel of Fig. 4. Here the low activity period is taken to be when the 10.7 cm radio flux is less than 80 s.f.u. while the high activity period is considered to be when the flux is more

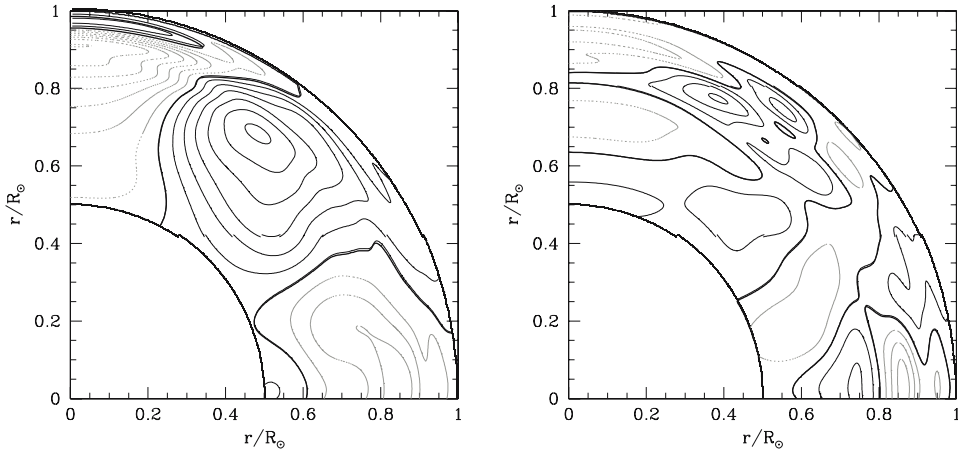


Figure 4. The contour diagram of the temporally averaged $\delta c^2 / c^2$. Because of symmetry only one quadrant is shown. The horizontal axis represents the equator, while the vertical axis is the rotation axis. The dark contours represent positive values while light contours show negative value. The thick contour marks the zero level. The contour spacing is 10^{-5} . The left panel shows the results using temporal averaging over all GONG datasets. The right panel shows the difference between the high-activity (>140 s.f.u.) and low-activity (<80 s.f.u.) periods.

than 140 s.f.u. There is a marginal variation in some parts which may be significant. Similar conclusions were obtained by Antia *et al.* (2003).

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Helioseismic Effects of Energetic Transients

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Abstract. Photospheric and chromospheric signatures related to large, energetic transients such as flares and CMEs, have been extensively reported during the last several years. In addition, energetic solar transients are expected to cause helioseismic effects. Some of the recent results are reviewed here; in particular, the helioseismic effects of the powerful flares in superactive region, NOAA 10486, including the 4B/X17 superflare of October 28, 2003. We also examine the temporal variations of power in low- l modes during the period May 1995–October 2005, and compare with daily, disk-integrated flare- and CME-indices to infer the effect of transients on the scale of whole solar disk.

Key words. Helioseismic effects—solar activity—solar transients.

1. Introduction

Solar flares release large amounts of energy at different layers of the solar atmosphere, including at the photosphere in the case of exceptionally major events. Therefore, it is expected that large flares would be able to excite acoustic waves on the solar surface, thereby affecting the p -mode oscillation characteristics. Acoustic modes that are always present on the Sun are generally accepted to be excited by turbulence in the convection zone (e.g., Goldreich & Kumar 1988). It was first suggested by Wolff (1972) that flares could excite solar oscillations as a result of the mechanical impulse produced by the thermal expansion exerted by a large flare towards the solar interior. He estimated the damping times to be longer than a day for the free modes. However, the earlier attempts to detect flare associated effects were mostly contradictory and inconclusive. For example, Haber *et al.* (1988) found a 14% greater power in the flaring region, while Braun and Duvall (1990) found no such effect for the energetic flares in NOAA AR 5395 of March 10, 1989. The difficulty in detecting any flare-related change is caused by the absorption of mode power by large sunspots which can absorb as much as 50–70% of the power of the p -modes (*c.f.*, Abdelatif *et al.* 1986; Braun *et al.* 1987; Rajaguru *et al.* 2001). Therefore, any excitation induced by the shorter-lived flares has to essentially compete with the effects of absorption associated with the intense magnetic fields of the sunspots.

Ambastha *et al.* (2003) have investigated characteristics of high-degree p -modes in several active regions of Cycle 23 that produced large flares. Using ring diagram technique, they found that the power in high-degree p -modes was larger during the period

of high flare activity as compared to non-flaring regions of similar magnetic field. Several workers have also studied the large, energetic flares that occurred in NOAA 10486 during October–November 2003 (Ambastha *et al.* 2004; Donea & Lindsey 2005; Kosovichev 2006). It is expected that the most detectable changes should occur during the most energetic flares. In the following sections, we present a review of recent studies of the helioseismic response associated with the super-flare of October 28, 2003 using different techniques. Finally, we discuss the correlation of disk-averaged daily flare- and CME-indices with power in the low- l modes during Cycle 23.

2. The 4B/X17 super-flare of October 28, 2003

Observations of photospheric magnetic fields have shown abrupt and persistent changes in time scales of the order of 10–100 minutes during flares. These are much larger than the time scales of a few minutes associated with the expected photospheric line profile changes with transients. In order to distinguish between transients, normal field evolution, and the abrupt and persistent changes in the magnetic field to be associated with solar flares, a few hour long data-cube of high magnetic sensitivity, high-cadence, high spatial resolution data are needed. Similarly, for studies pertaining to helioseismic response, a day long data-cubes of dopplergrams are required before, during and after the flare. GONG+ and SOHO-MDI provide the required data to explore flare-associated effects.

High cadence H_α filtergrams were obtained from Udaipur Solar Observatory (USO) providing chromospheric coverage of the active region NOAA 10486 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 the photospheric doppler velocity, line-of-sight magnetic field and the chromospheric H_α observations, the sites of velocity flows, magnetic flux emergence, and chromospheric restructuring in the active region were identified.

From the GONG+ magnetogram images obtained during the super-flare of October 28, 2003, “magnetic” features were observed moving rapidly from the flare site in a short duration from 10:56 UT to 11:15 UT. Both GONG and MDI data showed this feature. A portion of the feature was observed moving away from the neutral line, near the site of the flare onset towards the leading sunspot, and the other perpendicularly to it (Fig. 1). The scenario appeared like a blast wave, or Moreton wave observed in chromosphere (Fig. 2). Velocity of the moving magnetic feature was found to be of the order of 40 km/s, i.e., much larger as compared to the usual separation speed of two-ribbon flares, but similar to the velocity ≈ 45 km/s of seismic waves reported

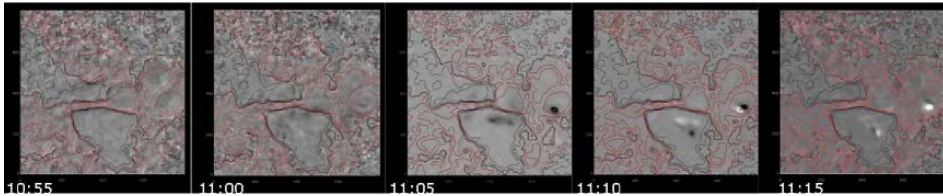


Figure 1. A moving “magnetic” feature propagating away from the site of the 4B/X17 flare as seen in the time-sequence of successive difference GONG magnetogram images.

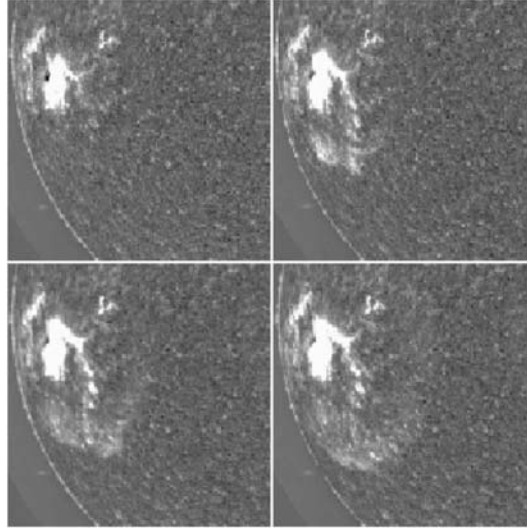


Figure 2. Moreton wave from the site of a X6 flare of 6 December 2006.

by Kosovichev and Zharkova (1998). If the moving feature is caused due to line profile changes, it is expected to show a relationship with the slower speed of flare ribbons separation. Also, it does not appear to be instrumental artifact either, as both the GONG+ and MDI images exhibit this feature. It is to note that GONG magnetic field measurements are not found very sensitive to line shape variation as shown by computer simulation of flare (Edelman *et al.* 2004). The exact nature of the moving feature requires further detailed investigation.

3. Helioseismic response of large flares

Locally excited acoustic modes interact with local variations in sound speed, flow fields, and magnetic activity which alter their propagation characteristics. A careful analysis of the acoustic waves in a localized patch of the solar photosphere can potentially reveal a great deal about the subsurface dynamics. Local helioseismology investigates small-wavelength acoustic waves that are confined principally to the near-surface layers, i.e., $r \leq 0.97 R_{\odot}$. There are several local helioseismology techniques in use, including the ring-diagram analysis (Hill 1988), time-distance method (Kosovichev *et al.* 2000), and acoustic holography (Lindsey & Braun 2000). Some of the important results derived from these are mapping of horizontal flows in the surface layers of the Sun revealing the meridional and zonal circulation patterns, as well as, smaller-scale flows associated with active regions and supergranulation, or solar subsurface weather, (SSW). Local helioseismology has also been used to study the acoustic and flow structure underlying sunspots, and to image active regions on the far side of the Sun (*c.f.*, Gizon & Birch 2005, for a review).

3.1 *Sun-quakes associated with flares – A photospheric wave phenomena*

Response of flares in solar atmosphere is observed in a variety of effects such as Moreton waves, eruptive mass ejections, X-ray, $H\alpha$ and white-light brightenings,

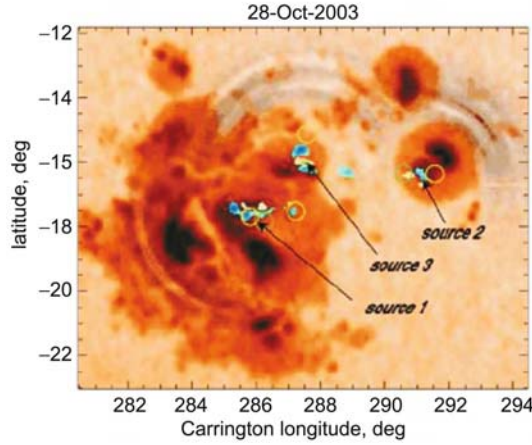


Figure 3. A white-light image of active region NOAA 10486 observed on October 28, 2003, superimposed with the Doppler signal at the impulsive phase, 11:06 UT (blue and yellow spots show up and down photospheric motions with variations in the MDI signal stronger than 1 km s^{-1}), positions of three wave fronts at 11:37 UT, and also locations of the hard X-ray (50–100 keV) sources (yellow circles) at 11:06 UT and 2.2 MeV γ -ray sources (green circles) (adopted from Kosovichev 2006).

etc. Waves have been observed everywhere on the Sun including the helioseismic, photoseismic, chromoseismic, as well as corona-seismic phenomena.

Traditionally, traveling waves or *Moreton waves* have been observed to originate from some flaring sites in the chromosphere (Fig. 2). A theoretical model proposed by Kosovichev & Zharkova (1995) predicted traveling waves or seismic waves, as opposed to the standing waves which constitute the normal modes of solar oscillations. These seismic waves are predicted to propagate away from the site of energetic flares with speeds increasing with distance. Kosovichev & Zharkova (1998) later observed the predicted seismic waves associated with a X2.6 flare of July 9, 1996. The source of the seismic response was suggested to be strong shock-like compression wave propagating downwards in the photosphere (Kosovichev 1986).

Observations of the seismic waves generated by the sunquake of 9 July 1996 is remarkably similar to the pattern of ripples generated by a stone dropped into a pond (Podesta 2003). They used a model of an inviscid, incompressible fluid in analogy to water waves, and found that the distances between successive wave crests were larger than observed, and concluded that the sunquake is composed primarily of acoustic (p -mode) waves rather than f -modes. Detection of sunquakes depends on their amplitude relative to the background solar noise. Presumably, all flares are expected to have seismic consequences at some level, but because of the high solar noise, the seismic waves may not easily be seen on individual Dopplergram images. It is easier to recognize these as expanding wave fronts in Dopplergram movies.

Seismic wave-like feature for the X17 superflare of October 28, 2003 in NOAA 10486 has been recently reported by Kosovichev (2006) using SOHO-MDI data. This was one of the largest events observed ever. It is interesting to note that some flares, having much weaker soft X-ray emission, produced higher amplitude seismic waves than the X17 flare (Kosovichev 2006). The study of sunquakes revealed some interesting features: (i) highly anisotropic nature of the seismic waves; (ii) the delay

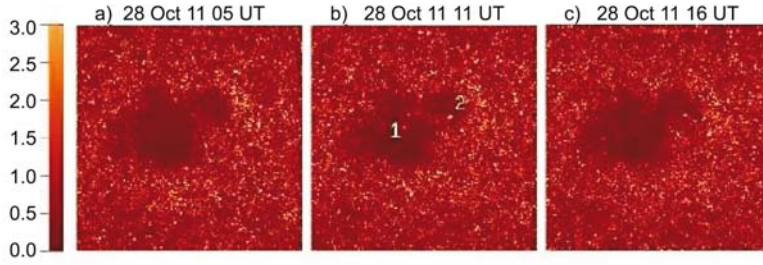


Figure 4. Egression power images for flare of October 28, 2003 obtained before, during and after the flare, showing compact acoustic sources of seismic waves (adopted from Donea & Lindsey 2005).

of 15–20 minutes after the initial impact, reaching the highest amplitude 20–30 minutes after the flare; (iii) traveling to large distances, exceeding 120 Mm, but in some cases, decaying more rapidly; (iv) propagation through sunspots without much distortion or significant decay, thus showing no evidence for conversion into other types of MHD waves. These observations indicate that the wave absorption by magnetic field is perhaps much weaker than expected by previous studies. The weak interaction of the seismic waves with sunspots can be explained by the fact that acoustic waves on the Sun do not propagate horizontally along the surface but travel through the deep interior. Therefore, the acoustic waves pass through the subsurface sunspot regions where the magnetic pressure is much smaller than the gas pressure, and reappear at the surface without much effect. The penetration depth of these waves depends on the travel distance. Therefore, the magnetic effects perhaps play a role only for shorter distances (i.e., smaller depths). The interaction of sunquakes with sunspots requires further analysis as the observations and analysis of sunquake events opens opportunities for developing new methods of helioseismic analysis of flaring active regions, similar to the methods of earthquake seismology.

3.2 Seismic sources of *p*-mode power associated with flares

Using helioseismic holography to image the seismic sources of the waves, Donea *et al.* (1999) obtained “egression” power maps for high-frequency acoustic waves during the flare studied by Kosovichev and Zharkova (1998). Interestingly, sunquake signals were not reported for any subsequent solar flares, thereby implying that sunquakes are rather rare phenomenon. However, recently Donea & Lindsey (2005) have reported that several flares in the declining phase of the Cycle 23 did show strong “egression” signals, which may be indicative of potential sunquakes. They have also detected compact sources of seismic waves emitted from powerful solar flares that occurred in NOAA 10486 on 2003 October 28 and 29 (Fig. 4). Interestingly, these acoustic sources were strongly associated with the foot-points of a coronal loop that hosted the flares and hard X-ray signatures.

3.3 Amplification of *p*-mode power by flares in NOAA 10486

The effect of flares on solar oscillation modes has been explored for several active regions using ring diagram analysis (Ambastha *et al.* 2003). This study gives evidence

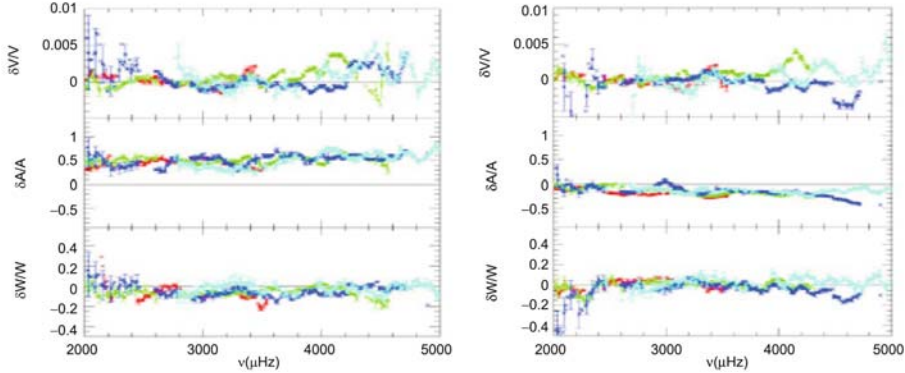


Figure 5. Relative difference in mode frequencies, widths and ratio of peak power for NOAA 10486 using ring-diagram analysis. The left and right panels respectively show the results for p -modes with $n = 0, 1, 2, 3$ obtained for the flare/preflare (i.e., Oct 28/Oct 27) and the flare/postflare (i.e., Oct 28/Oct 29) phases (adopted from Ambastha *et al.* 2004).

of flare-associated variation in mode characteristics such as frequency, width, power and asymmetry parameter. Average variation in p -mode amplitude for 11 active regions showed a good correlation with flare index after correcting the mode amplitude for magnetic field strength of the active region (*cf.*, Rajaguru *et al.* 2001). For the 4B/X17 superflare of October 28, 2003 also it was found that the p -mode power increased substantially after the flare. In fact, the mode-power increased even further on October 29, 2003 as a result of more flares occurring in the active region as shown in Fig. 5. Recent studies have shown that subsurface flows may also lead to p -mode characteristic variations (Zhao & Kosovichev 2004; Haber *et al.* 2004).

4. Subsurface signatures of flaring active regions

There have been searches for the subsurface signatures of large flares. For example, it is interesting to compare the velocity profiles with depth in flare-active and flare-inactive regions (Ambastha *et al.* 2004). The meridional velocity, u , in flaring active regions such as NOAA 10486, is found to possess a steep gradient below a depth of 4 Mm (Fig. 6a). Similar behaviour was obtained for active regions NOAA 9026 and 9393 (Ambastha *et al.* 2003), that appears to be a common feature of flare-productive active regions. In some regions, e.g., NOAA 9026, it was found that the steep gradient vanished after the flare (Ambastha *et al.* 2004). Similarly, Komm *et al.* (2005) found strong signal in each of the vorticity components, and hence in the kinetic helicity as a function of latitude and depth at the location of the active region during the epoch when the flares occurred (Fig. 6b). The sign of the kinetic helicity remained the same at depths greater than about 5 Mm at the locations of AR 10486, while closer to the surface, the sign changed with depth indicating a more complicated behaviour. This is similar to the result obtained by Ambastha *et al.* (2004). Horizontal flows surrounding NOAA 10486 and variation with depth in the upper 14 Mm of the convection zone were studied by Haber *et al.* (2004). They reported large-scale shear flows that may contribute to conditions conducive to intense flaring activity.

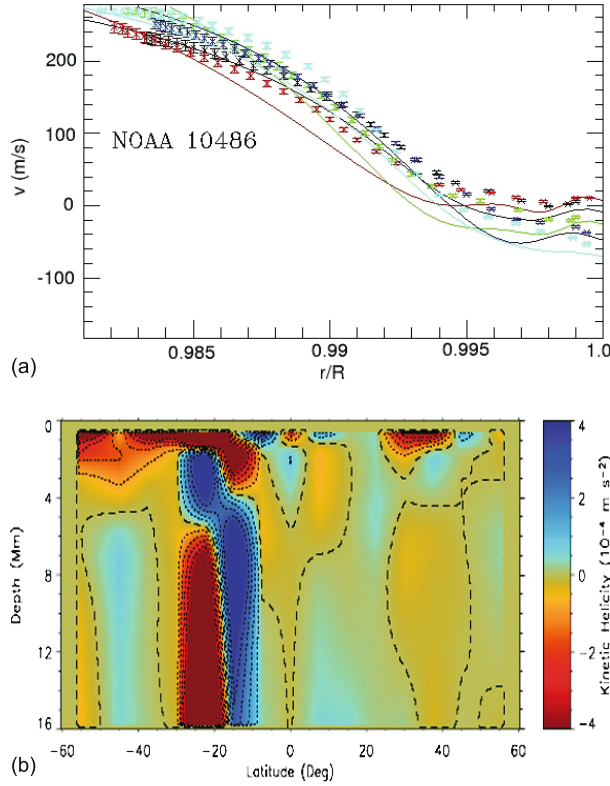


Figure 6. (a) The meridional (u) and zonal (v) components of velocity for NOAA 10486 at different stages of evolution. The *lines* show the results obtained by *RLS* inversions while *points* mark those obtained by *OLA* (adopted from Ambastha *et al.* 2004). (b) Kinetic helicity with depth. The sign of kinetic helicity remains the same at depths greater than about 5 Mm at the location of NOAA 10486. Closer to the surface, the sign changes with depth indicating a more complicated behaviour (adopted from Komm *et al.* 2005).

5. Excitation of p -modes by energetic transients on the global scale

The acoustic modes, always present on the Sun, are generally expected to be excited by turbulence in the convection zone (Goldreich and Kumar 1988). The energetic transient phenomena, that contribute additional energy to these modes at the local spatial scales of active regions, could also be expected to contribute energy to the solar modes on a global scale. For example, Chaplin *et al.* (1997) found excess of excitation events at the highest powers in comparison with the model based on stochastic excitation of modes (Goldreich *et al.* 1994). Gavryusev & Gavryuseva (1999) reported that the power distributions derived from radial mode observations contained a significant pulse of power at large energies, which they attributed to energetic coronal mass ejection (CME) events. However, Chaplin *et al.* (2004) found no convincing response of CMEs on the solar mean magnetic field data acquired by the BiSON Instruments using the correlation technique. The existence of high power events in the p -mode data, if indeed found to be significant, raises the question of their origin that needs to be addressed.

In order to examine whether disk-integrated flare and CME activity have such effects, Ambastha & Antia (2006) re-investigated the temporal variation of power in low- l modes using a 3816-day long time-series of mode power available from GONG over the period May 1995–October 2005, covering the solar minimum to maximum and a good part of the descending phase of Cycle 23. Theoretical model of mode excitation by turbulence predicts a random distribution of χ^2 -type for the running mean power, decreasing monotonically after maximum with increasing $E_T/\langle E \rangle$ (Kumar *et al.* 1988), where E_T represents the mode energy and $\langle E \rangle$ is the average mode energy over the interval of interest. According to the theoretical model, as T is increased, the width of the distribution function narrows, the peak shifts towards the mean energy, and eventually the distribution approaches a δ -function at very large T . A long-period time series of mode power, such as the time-series now available from the GONG network, is required for obtaining the observed distribution function and its comparison with the theoretical distribution. The comparison is expected to help resolve the problem of mode excitation by checking if there is a departure from the theoretically expected distribution at high power. Ambastha & Antia (2006) reported that the observed power distribution of low- l modes essentially followed the theoretically expected stochastic distribution with no significant features seen at large power for data averaged over large (or small) running windows. This finding supports the generally accepted mechanism of excitation of p -modes by turbulence. However, the mean level of running mean power was found to vary; decreasing from solar minimum to maximum, i.e., during the ascending phase and then increasing in the descending phase of solar Cycle 23. The temporal variation in power of the modes, between the minimum and maximum, was found to be about 20–30%; consistent with the known variation in mode power (e.g., Elsworth *et al.* 1993). Also, a mild anti-correlation is found on global scale with disk-integrated flare-index and CME-counts (Ambastha & Antia 2006).

6. Conclusions

Using ring diagram technique, significant increase in mode power of acoustic modes has been found during large flares, beyond the normal value expected from the influence of magnetic fields. Compact seismic sources of p -mode power were identified at some locations using holographic egression maps (Donea & Braun 2005). Propagating seismic waves were also observed from the site of the X17 superflare (Kosovichev 2006). Meridional velocity in the flaring region was found to possess a steep gradient below a depth of 5 Mm (Ambastha *et al.* 2004), and a similar feature was observed in kinetic helicity (Komm *et al.* 2005).

On the global scale of the solar disk, however, the observed power distribution of low- l modes followed the expected theoretical χ^2 -type distribution for mode excitation by turbulence. No feature at large power, corresponding to energetic transients, was found. In fact, the correlation of the daily, disk-averaged flare- or CME-indices with low- l mode power was found to be poor (Ambastha & Antia 2006). However, temporal variation in power of modes between the minimum and maximum of Cycle 24 was found to be about 20–30%, consistent with the known variation, e.g., Elsworth *et al.* (1993).

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Magnetic and Velocity Field Variations in the Active Regions NOAA 10486 and NOAA 10488

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Abstract. We study the magnetic and velocity field evolution in the two magnetically complex active regions NOAA 10486 and NOAA 10488 observed during October–November 2003. We have used the available data to examine net flux and Doppler velocity time profiles to identify changes associated with evolutionary and transient phenomena. In particular, we report detection of rapid moving features observed in NOAA 10486 during the maximum phase of the X17.2/4B superflare of October 28, 2003. The velocity of this moving feature is estimated around 40 km/s, i.e., much greater than the usual H α flare-ribbons' separation speed of 3–10 km/s, but similar to the velocity of seismic waves, i.e., ~ 45 km/s reported earlier by Kosovichev & Zharkova (1998).

Key words. Sun: magnetogram—Sun: Dopplergram—Sun: flare.

1. Introduction

During the declining phase of solar cycle 23, two large active regions NOAA 10486 and NOAA 10488 appeared in the period October 17–November 4, 2003. The active region NOAA 10486 was most active, producing flares of unprecedented magnitudes and fast moving coronal mass ejections. Rapid and permanent changes of photospheric magnetic shear associated with large flares were observed in earlier studies (Ambastha *et al.* 1993; Wang *et al.* 1994; Wang 2006). We report here rapid change in the photospheric magnetic flux in association with the large, energetic X17.2/4B flare observed in NOAA 10486 on October 28, 2003.

2. Magnetic and velocity field changes in the active regions

We found a correlation of $\approx 40\%$ in the net magnetic flux time-series data obtained for the two ARs, NOAA 10486 and NOAA 10488. SOHO-EIT image also clearly shows an interconnection between the two ARs through a trans-equatorial loop. We selected eight areas-of-interest, as marked in Fig. 1(a), along the neutral lines, forming narrow channels of strong magnetic field gradients around the location of flare ribbons, and in quiet regions away from the flare. We evaluated net magnetic flux over these boxes,

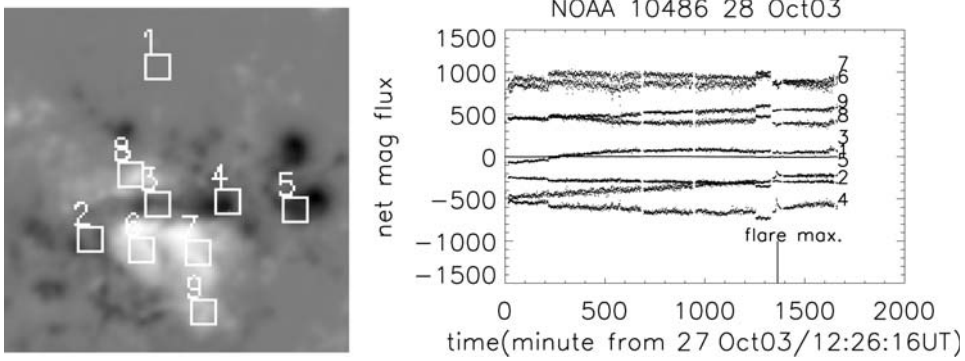


Figure 1. Temporal evolution of net magnetic flux in selected areas-of-interest (10×10 pixel size) in NOAA10486 during October 27–28, 2003.

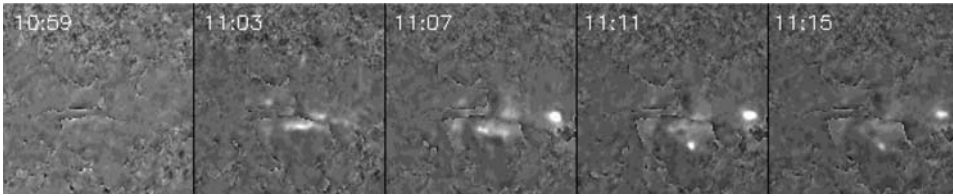


Figure 2. Evolution of magnetic flux and the moving feature observed during 10:59–11:15 UT seen in difference images ($I_t - I_{10:55}$).

10×10 pixel in size, during the day of flare, i.e., October 27–28, 2003 (Fig. 1b). Both gradual, evolutionary and abrupt, flare associated changes in magnetic flux are observed from the temporal profiles.

3. Moving “magnetic” features observed during the X17/4B superflare of October 28, 2003

From the GONG + magnetogram images obtained during the super-flare of October 28, 2003, a “magnetic” feature was observed moving rapidly from the flare site towards the leading sunspot in a short duration from 10:56 UT to 11:15 UT (Fig. 2). Both GONG and MDI data showed this feature. A portion of the feature was observed moving away from the neutral line, originating near the site of the flare onset, towards the leading sunspot and the other, perpendicularly to it. The scenario looks like a wave propagating away, like a blast wave, or Moreton waves, observed in chromosphere. Interestingly, no such feature was observed during the X10/3B flare on October 29, 2003, which also was a white light flare.

Using center-of-mass formula we found the velocity of the moving magnetic feature of the order of 40 km/s, which is much larger as compared to the usual separation speed of two-ribbon flares, but similar to the velocity ~ 45 km/s of seismic waves reported earlier by Kosovichev and Zharkova (1998). If the moving feature is caused due to line profile changes, it is expected to show a relationship with the speed of flare ribbons’ separation, which is usually much slower. Also, it does not appear to be instrumental

artifact either, as both the GONG + and MDI images exhibit this feature. It is to be noted that GONG magnetic field measurements are not found very sensitive to line shape variation as shown by computer simulation of flare models (Edelman *et al.* 2004). The exact nature of the moving feature would require further detailed investigation.

4. Conclusions

- A correlation ($\approx 40\%$) between the net magnetic flux time profiles of NOAA 10486 and 10488 was found on October 28, 2003, suggesting an interconnection between these ARs, which is further corroborated by SOHO-EIT image showing their interconnection by a trans-equatorial loop.
- Both gradual evolutionary and abrupt flare associated changes in the magnetic flux are observed in several locations in NOAA 10486 during the X17/4B superflare of October 28, 2003.
- A rapidly moving feature was observed originating from the flare site. It appeared like a wave propagating as Moreton wave observed in chromosphere. Its estimated speed ~ 40 km/s is much larger compared to the usual separation speed of 3–10 km/s of two-ribbon flares, but similar to that of the seismic waves ~ 45 km/s reported earlier by Kosovichev & Zharkova (1998). Exact nature of this feature needs further investigation.

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Software for Interactively Visualizing Solar Vector Magnetograms of Udaipur Solar Observatory

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Abstract. The Solar Vector Magnetograph (SVM) at Udaipur Solar Observatory saw its first light in April 2005. The retrieval of vector fields from the imaging spectro-polarimetric observational data requires a substantial amount of computer post-processing. The GUI-based data reduction and analysis software have been developed to make the data processing pipeline user-friendly and less time-consuming. In this paper we describe these software packages.

Key words. Instrumentation—polarimetry—spectroscopy.

1. Introduction

The magnetic field vector as measured by solar magnetographs can be displayed in more ways than one. The visualization of vectors is crucial to understanding the changes in field morphology. It is crucial to the transforming of vectors in heliographic co-ordinates and to applying corrections by performing geometric transformation for active regions especially when they are observed away from the disk center (Venkatakrishnan *et al.* 1988). These transformations are necessary before retrieving physical quantities from the observed magnetic field vectors like magnetic shear, etc.

2. The instrument

The schematic of the instrument is given in Fig. 1. The details about the instrument can be found elsewhere (Gosain *et al.* 2004, 2006). The instrument is basically a tunable Fabry–Perot-based imaging magnetograph which uses quarter wave-plates as polarization modulator and crossed calcites (Savart plate) as dual beam analyzer. The simultaneous measurements of orthogonal polarization components reduce the seeing induced spurious polarization signals. The straight and symmetric telescope design is used to avoid instrumental polarization to a great extent. The polarized profiles with single camera exposure can be obtained with rms noise level of $< 0.5\%$ in the continuum. The complete Stokes profile with 21 samples across the Fe I 6302.5 Å line can be performed in less than a minute with polarization sensitivity of $< 0.5\%$ of continuum.

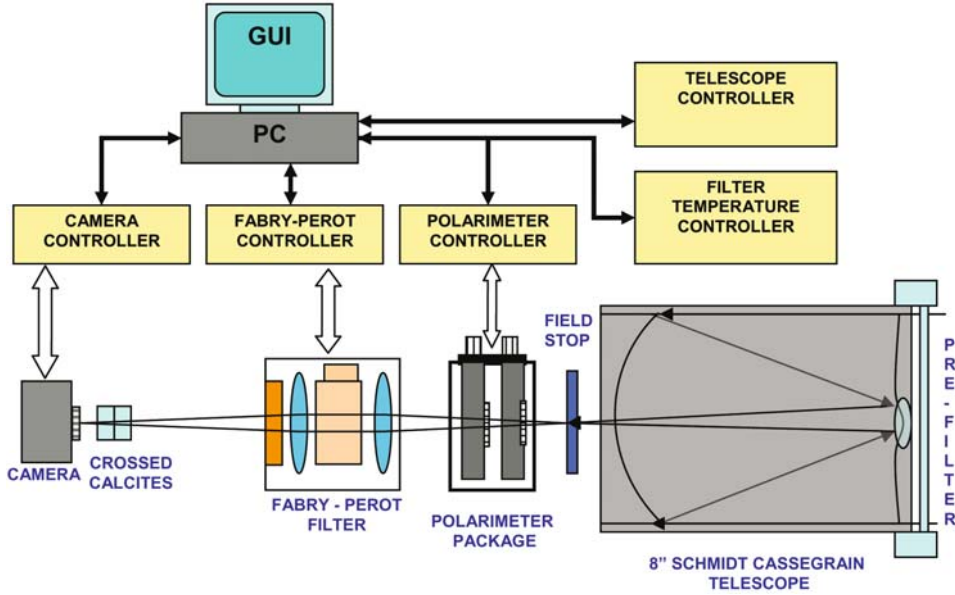


Figure 1. Schematic of Solar Vector Magnetograph (SVM) at Udaipur Solar Observatory.

3. Tools for data analysis and visualization

In order to facilitate data analysis and visualization of the calibrated data in a user-friendly way, we have developed two graphical user interfaces (GUIs). The first one is developed for performing interactive Stokes inversion of the observed profiles using a Milne–Eddington inversion scheme. While another GUI is developed for visualizing the magnetic field parameters retrieved by Stokes inversion. In the following sections we describe these software packages.

3.1 Stokes inversion

A GUI is developed for interactively inverting the Stokes profiles over the observed region (see Fig. 2) under Milne–Eddington atmosphere assumptions (Hector Socas-Navarro 2001). However, the complete inversion of the observed region is highly time consuming, taking about an hour for single dataset. A hybrid approach is being developed where weak field approximation (WFA) (Jefferies *et al.* 1991), is used for those pixels where the polarization signals are below a particular threshold. This threshold corresponds to the limit up to which WFA is valid. The threshold value is determined using synthetic Stokes profiles generated by MELANIE, whereas for pixels above this threshold, Stokes inversion is used to retrieve the magnetic field. This approach will save a lot of computational time since WFA is generally valid for a major portion of the observed area.

3.2 Magnetogram visualization

In order to visualize vector magnetogram a GUI has been developed (see Fig. 3). The inverted data can be visualized in the form of arrows showing transverse field

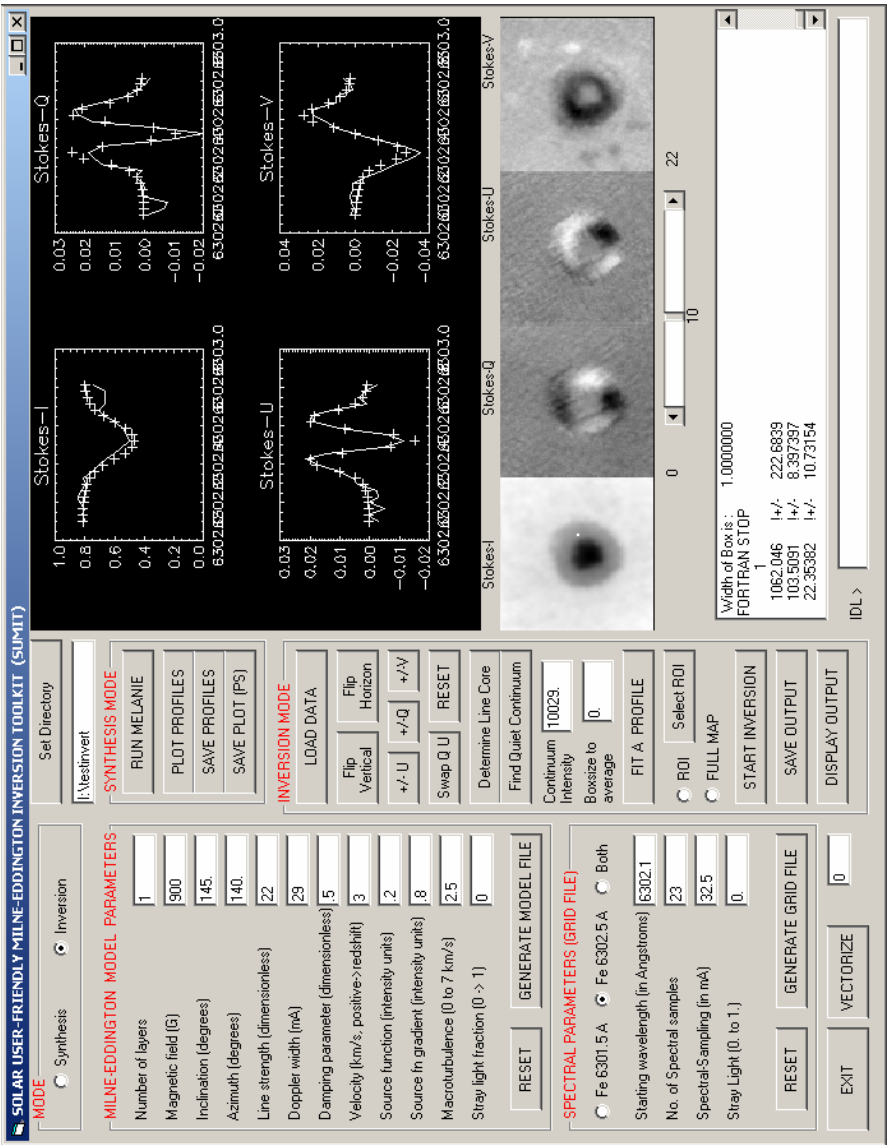


Figure 2. A graphical user interface for Milne–Eddington Inversion of SVM Stokes profiles.

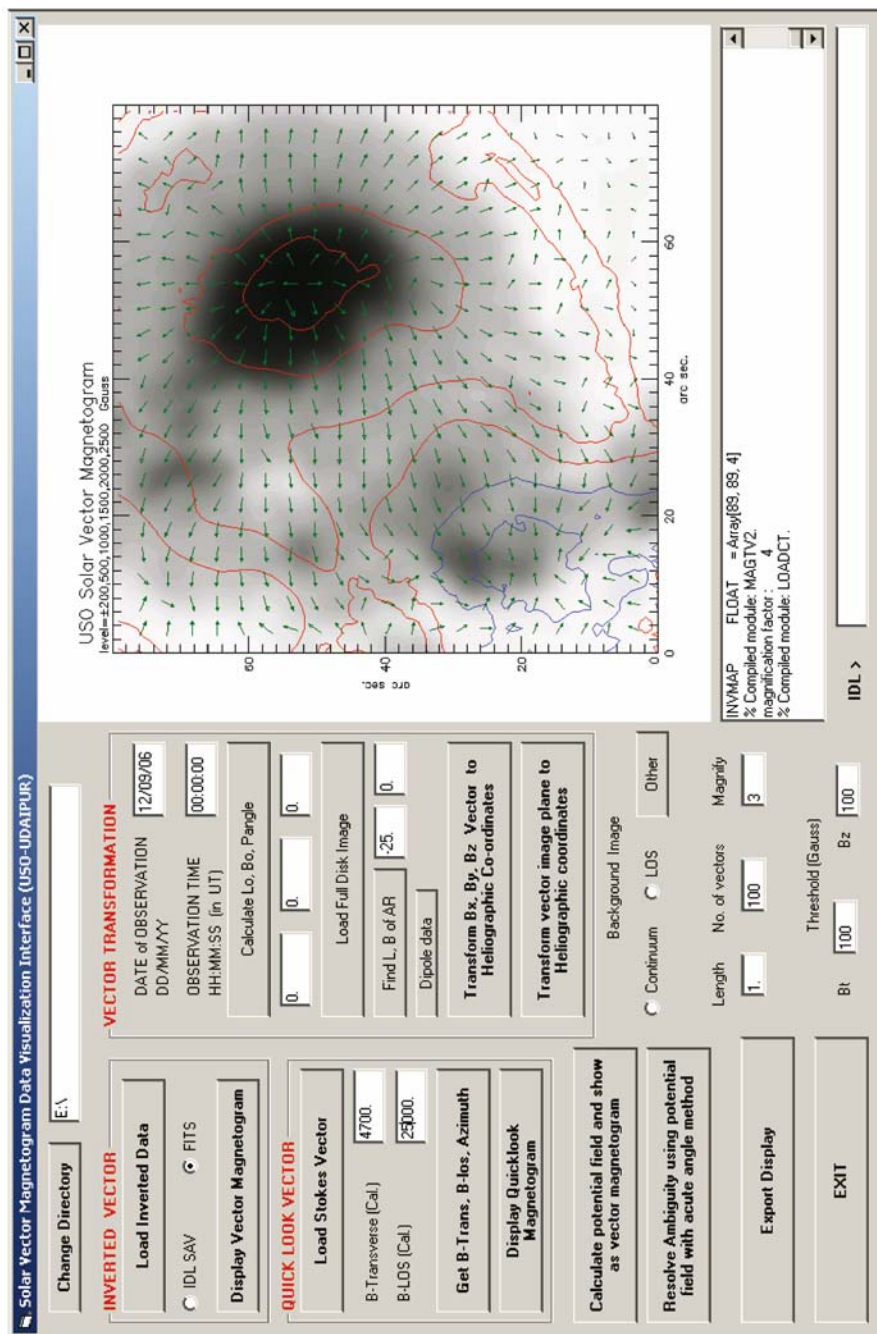


Figure 3. A graphical user interface for visualizing and transforming the inverted vector field data.

strength and direction while contours show line-of-sight magnetic field. The quick-look feature is also available for retrieving magnetic field vectors from polarized images in the line wing after determining suitable calibration constants. Further, the heliographic transformation of the vector magnetograms observed away from disk center can be performed interactively. The potential field can be calculated and displayed interactively. Also, the 180° ambiguity resolution can be performed using the acute angle method. The interface also provides utilities for calculation of L0, B0 and position angle for a given data and time. Also, one can load full disk images of the Sun and determine the L and B coordinates of active regions interactively by clicking on them.

4. Conclusions

These tools will prove to be helpful in future for a novice observer or user of the instrument to perform data analysis and generate its visualizations. Further, these tools are also useful for retrieving the magnetic field parameters at particular locations of interest without inverting the entire maps. Also, it will allow users to change the guess parameters of the models from the interface itself and there will be no need to edit the parameter file manually. These interfaces are quite general in their design and can accept any other solar observations as input with little reformatting. Thus, these tools will prove useful to other users of the solar community. Also, for new instruments developed henceforth at Udaipur Solar Observatory producing spectro-polarimetric data, these tools can be used with little reformatting in the input data. As, these GUIs are based on Visual BasicTM and IDL software, we can modify or upgrade the softwares by just replacing or adding new IDL routines.

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Solar Flare Physics Enlivened by TRACE and RHESSI

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Abstract. The Transition Region and Coronal Explorer (TRACE) gave us the highest EUV spatial resolution and the Ramaty High Energy Solar Spectrometric Imager (RHESSI) gave us the highest hard X-ray and gamma-ray spectral resolution to study solar flares. We review a number of recent highlights obtained from both missions that either enhance or challenge our physical understanding of solar flares, such as:

1. Multi-thermal Diagnostic of 6.7 and 8.0 keV Fe and Ni lines,
2. Multi-thermal Conduction Cooling Delays,
3. Chromospheric Altitude of Hard X-Ray Emission,
4. Evidence for Dipolar Reconnection Current Sheets,
5. Footpoint Motion and Reconnection Rate,
6. Evidence for Tripolar Magnetic Reconnection,
7. Displaced Electron and Ion Acceleration Sources.

Key words. Solar flares—particle acceleration—hard X-rays.

1. Introduction

In this review, we touch on some selected highlights in the recent exploration of solar flare physics, mostly obtained from data of the TRACE and RHESSI missions. These two missions are most instrumental in the progress of our physical understanding of solar flares. At the time of writing we count some 690 TRACE publications since its launch in 1998, and some 330 RHESSI publications since its launch in 2002. RHESSI provides us images and spectra of the bremsstrahlung of high-energy particles in flares, while TRACE provides us EUV images of the coronal post-flare loop system once it cools down to $\lesssim 2$ MK, as well as from plasma at 15–20 MK (through the Fe XXIV line at 195 Å) and thus reveals the context of the magnetic configuration. In addition, UV images from TRACE show us also the chromospheric flare ribbons, which have an intimate connection to the hard X-ray footpoint sources.

RHESSI is a Fourier imager of the class of rotation-modulated collimators (RMC), with 9 bi-grid collimators that have grid pitches from 2.3'' to 180'', each one with a sensitive area of 39 cm². The rotation period of the spinning spacecraft is maintained at ≈ 4 s. Cooled germanium detectors provide an energy resolution with a FWHM of $\lesssim 1.0$ keV. Imaging is accomplished in the energy range of $\gtrsim 3$ keV

up to gamma rays $\lesssim 17$ MeV. RHESSI was launched on 2002 February 5 and detected more than 20,000 flares during the 2002–2006 period of the declining solar cycle 23.

TRACE has a 30-cm aperture telescope with normal-incidence coatings for EUV and UV in 4 quadrants. An internal pointing stabilizer limits the spacecraft jitter to $\lesssim 0.1''$. Images are read out from a CCD detector with a size of 1024×1024 and a pixel size of $0.5''$ (360 km on the Sun). The field-of-view is $8.5''$, roughly a quarter of a solar diameter. The 4 EUV channels have the wavelengths of 171 Å (Fe IX, Fe X), 195 Å (Fe XII, Fe XXIV), 284 Å (Fe XV), and 304 Å (He II). In addition TRACE has UV channels at 1216 Å (H I, Ly α), 1550 Å (C IV), 1600 Å, and 5000 Å (white-light). TRACE was launched in 1998 April and recorded several million EUV and UV images since.

2. Multi-thermal diagnostic of 6.7 and 8.0 keV Fe and Ni lines

A typical energy spectrum of a large solar flare is dominated by thermal (free-bound) bremsstrahlung at $\epsilon \approx 1$ –20 keV (soft X-rays), collisional bremsstrahlung of non-thermal electrons at $\epsilon \approx 20$ –1000 keV (hard X-rays), nuclear de-excitation lines at $\epsilon \approx 1$ –10 MeV (gamma-ray lines), collisional bremsstrahlung of high-energy electrons at ≈ 10 –100 MeV, and pion decay radiation at ≈ 0.2 –5 GeV. While the thermal spectrum at $\epsilon \lesssim 20$ keV mostly showed a featureless exponential-like shape with previous hard X-ray detectors (e.g., SMM, CGRO, Yohkoh/HXT), the new germanium-cooled detectors of RHESSI (Lin *et al.* 2002) and the solid-state detectors of the Soft X-ray Spectrometer (SOXS) (Jain *et al.* 2005) onboard the *GSAT2* Indian spacecraft provide sufficient high energy resolution to resolve the 6.7 and 8.0 keV Fe and Fe/Ni line complexes (Phillips 2004; Phillips *et al.* 2006). An example of a flare spectrum with resolved 6.7 keV Fe and 8.0 keV Fe/Ni lines resolved during a flare is shown in Fig. 1, recorded by SOXS.

The 6.7 keV Fe line is detectable for flare plasma temperatures in the range of $T \approx 8$ –35 MK and provides a new diagnostic on the flare temperature. The predicted equivalent width of the 6.7 keV line complex steadily increases from zero at $T \approx 8$ MK up to ≈ 3 keV at $T \gtrsim 20$ MK (Fig. 2 in Dennis *et al.* 2007; Phillips *et al.* 2006). It was observed to exceed the theoretically expected one in the temperature range of $T \approx 20$ –30 MK (Fig. 2 in Dennis *et al.* 2007), which could indicate multi-thermal flare plasma or a higher iron abundance. The ratio of the flux in the 6.7 keV complex to the flux in the 8.0 keV complex is expected to decrease progressively with higher plasma temperature, which has been observationally confirmed for the 2002 July 23 flare (Fig. 3 in Dennis *et al.* 2007). Thus, these two-line complexes provide sensitive new diagnostic of the flare plasma temperature and iron abundance.

3. Multi-thermal conduction cooling delays

Since major solar flares generally produce a large number of individual postflare loops, giving the familiar appearance of loop arcades lined up along the flare ribbons, it is unavoidable that each loop is heated and cools off at different times so that a spatially integrated spectrum always contains a multi-thermal differential emission measure

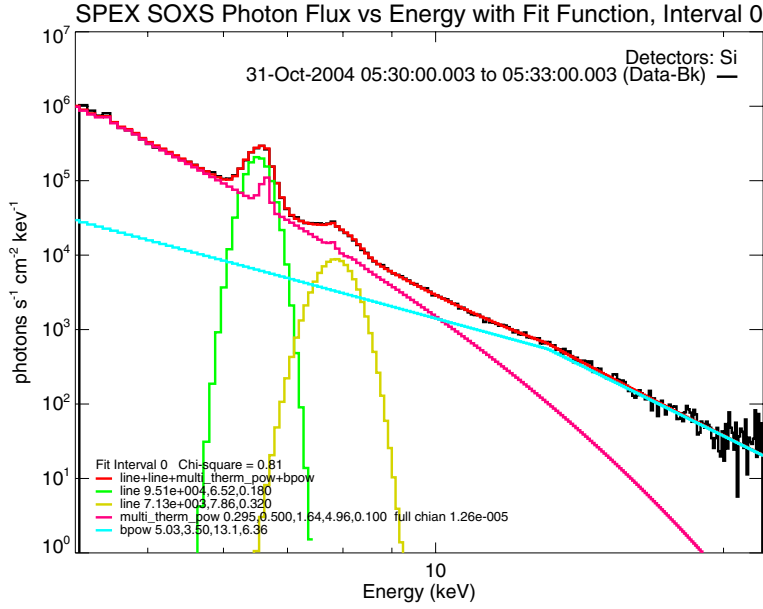


Figure 1. Photon flux spectrum of the 2004 October 31, 05:30 UT, flare, observed by SOXS. The thermal spectrum is fitted with a multi-thermal (exponential-like) spectrum plus a nonthermal (Powerlaw) spectrum, as well as with two gaussians for the 6.7 keV Fe and 8.0 keV Fe/Ni lines. The spectral resolution of SOXS is ≈ 0.7 keV. The nonthermal component may contain pulse pile-up events, based on comparisons with RHESSI spectra (courtesy of Rajmal Jain).

distribution. The resulting multi-thermal bremsstrahlung spectrum (for photon energies ϵ) observed in soft X-rays (neglecting the Gaunt factor of order unity),

$$F(\epsilon) = F_0 \int \frac{\exp(-\epsilon/k_B T)}{T^{1/2}} \frac{dEM(T)}{dT} dT, \quad (1)$$

is then a function of a multi-thermal differential emission measure distribution $dEM(T) = n^2(T)dV$. The initial cooling of the hot flare plasma (say at $T \gtrsim 10$ MK) is generally dominated by conductive cooling (rather than by radiative cooling, which dominates later on at EUV temperatures of $T \lesssim 2$ MK). The thermal conduction time has the following temperature dependence:

$$\tau_{\text{cond}}(T) = \frac{\epsilon_{\text{th}}}{dE/dt_{\text{cond}}} = \frac{3n_e k_B T}{d/ds \kappa T^{5/2} dT/ds} \approx \frac{21}{2} \frac{n_e L^2 k_B}{\kappa} T^{-5/2}. \quad (2)$$

Since thermal bremsstrahlung involves an increasing photon energy with higher flare plasma temperature, the conductive cooling time is expected to become shorter with higher temperatures ($\tau_{\text{cond}} \propto T^{-5/2}$), or vice versa, longer cooling delays are expected at lower energies. This delay is related to the Neupert effect (Dennis & Zarro 1993), which predicts that the soft X-ray light curve mimics the integral of the hard X-ray light curve. Thus, the soft X-ray peak is always delayed with respect to the hard X-ray peak, essentially reflecting the time delay between chromospheric heating, evaporation, and cooling of flare loops. This conductive cooling delay $\tau_{\text{cond}}(\epsilon)$ has been measured and

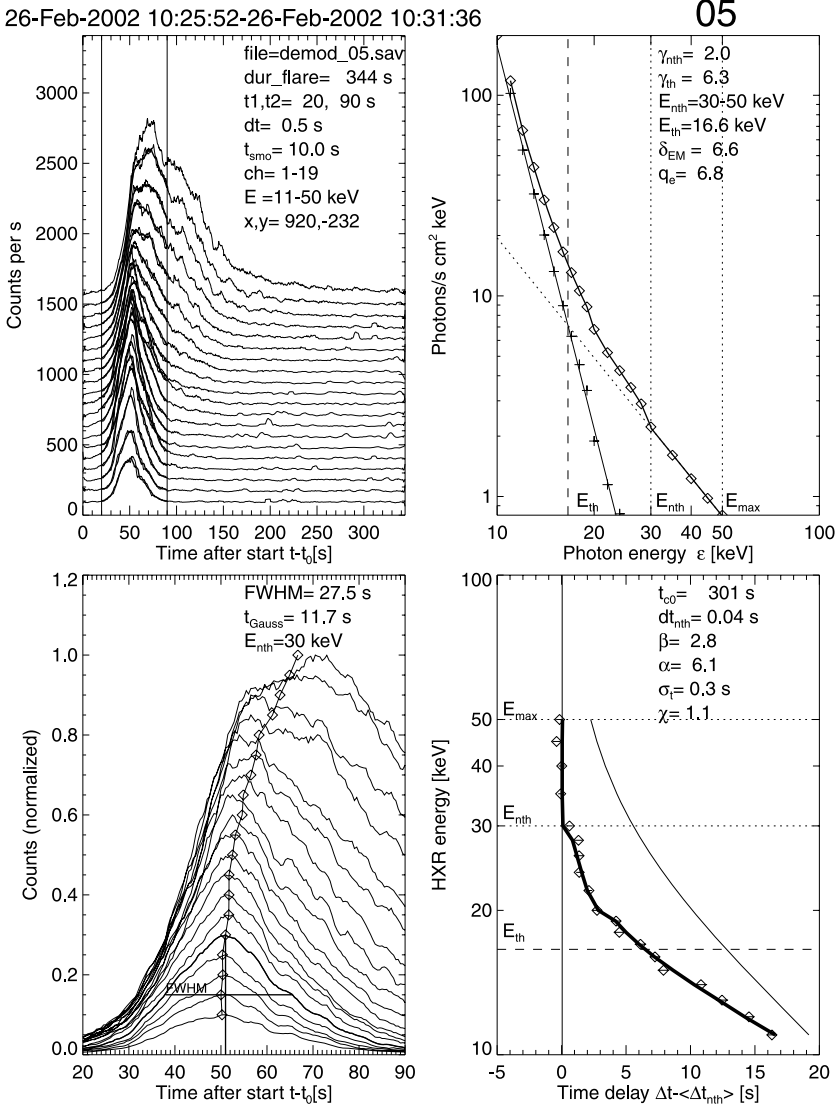


Figure 2. X-ray light curves are shown for the 2002 February 26, 10:31 UT, flare, for energies of 10 keV to 30 keV in intervals of 1 keV, observed with RHESSI (left panels). The spectrum is decomposed into thermal and nonthermal components (top right panel) and the delay of the peaks at different energies is fitted with a thermal conduction cooling time model that has a scaling of $\tau_{cond}(T) \approx T^{-\beta}$ (right bottom panel). The best fit shows a power index of $\beta = 2.8$, which is close to the theoretically expected value of $\beta = 5/2$ (equation 2). The full delay of the thermal component is indicated with a thin curve (bottom left panel), while the weighted (thermal + nonthermal) fit is indicated with a thick curve (from Aschwanden 2007).

modeled for a comprehensive set of RHESSI-observed flares (see example in Fig. 2), and yields a new diagnostic of the process of conductive cooling in multi-thermal flare plasmas (Aschwanden 2007).

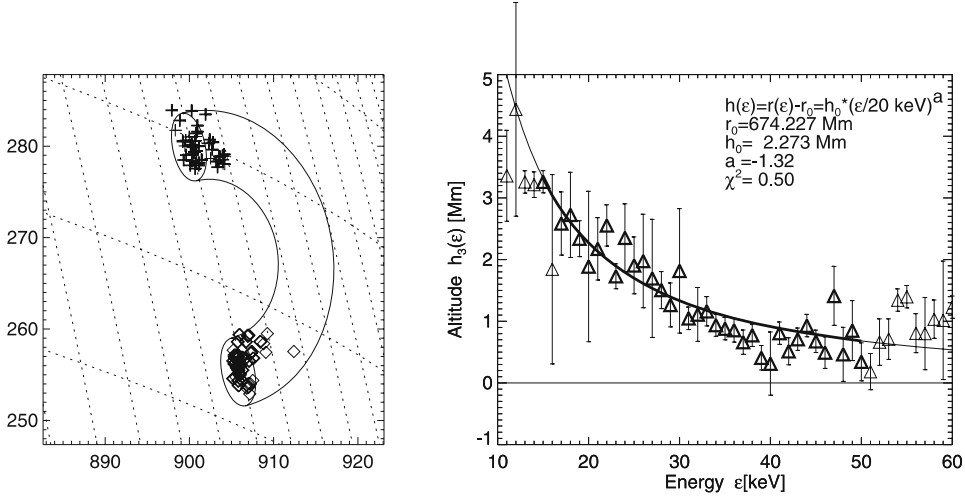


Figure 3. The centroids of footpoint hard X-ray emission are marked for different photon energies between 10 keV and 60 keV for the 2002 February 20, 11:06 UT, flare, which occurred near the solar west limb and was imaged with RHESSI (left panel). The altitude $h(\epsilon)$ as a function of energy ϵ shows a systematic height decrease with increasing energy (right panel) (from Aschwanden *et al.* 2002).

4. Chromospheric altitude of hard X-ray emission

Hard X-ray footpoint sources result from collisional bremsstrahlung of precipitating electrons, which produce most of the emission in the chromosphere according to the thick-target model. Essentially, mildly energetic electrons lose their energy already in the lower corona or transition region, while the more energetic high-relativistic electrons penetrate deeper into the chromosphere. The altitude of these hard X-ray footpoint sources could never be measured accurately before RHESSI, because of a lack of spatial and spectral resolution. With RHESSI we can measure the centroid of the footpoint location within an accuracy of order an arcsecond for every photon energy in steps of 1 keV. For a flare near the limb (Fig. 3), the centroid location translates directly into an altitude. Such a flare was observed on 2002 February 20, where the heights of the footpoint sources were fitted with an exponential function of the photon energy, which yielded altitudes in the range of $h \approx 1000\text{--}5000$ km in the energy range of $\epsilon = 10\text{--}60$ keV, progressively lower with higher energy, as expected from the thick-target model (Fig. 3, right frame). Since the stopping depth of the precipitating electrons is a function of the integrated density along their path in the chromosphere, the measured height dependence of the hard X-ray centroids can be inverted to yield a density model of the chromosphere. The inversion of the RHESSI data in the example shown in Fig. 3 yields a chromospheric density model that has a significantly higher electron density in the $h = 2000\text{--}5000$ km range (Aschwanden *et al.* 2002) than the standard chromospheric models based on UV spectroscopy and hydrostatic equilibrium (VAL and FAL models). The RHESSI-based chromospheric density model is therefore more consistent with the “spicular extended chromosphere”, similar to the sub-mm radio observations during solar eclipses carried out at CALTECH (Ewell *et al.* 1993).

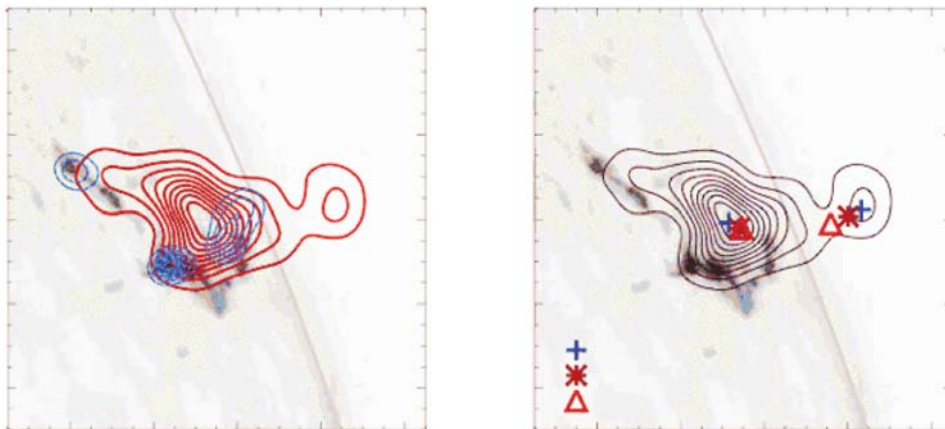


Figure 4. **Left:** TRACE 1600 Å image of the 2002 April 15, 23:11 UT, flare, overlaid with RHESSI contours of thermal and nonthermal emission. **Right:** Same image with contours of RHESSI 10–15 keV emission. The symbols indicate the centroid of the coronal sources in the energy bands of 6–8 keV, 10–12 keV, and 16–20 keV for the lower coronal source, and 10–12 keV, 12–14 keV, and 14–16 keV for the upper coronal source. Note the increasing separation of the coronal hard X-ray sources towards lower thermal energies (from Sui & Holman 2003).

5. Evidence for dipolar reconnection current sheets

The standard (Carmichael–Sturrock–Hirayama–Kopp–Pneuman) flare model envisions two oppositely directed magnetic field regions that are stretched in the vertical direction to form a current sheet where they reconnect. After reconnection, the newly-connected field lines form a cusp beneath the X-type reconnection point and relax into a semi-circular, dipolar post-flare loop, which has two conjugate footpoints where the nonthermal hard X-ray emission originates. In a recent unique RHESSI observation (Sui & Holman 2003), two coronal hard X-ray sources were observed, symmetrically placed above and below the putative X-point, at locations where the downward and upward outflows from the reconnection region are expected (Fig. 4). Moreover, the separation of the sources increased with lower energies in the thermal range of ≈ 10 –16 keV. Since increasing temperatures affect higher photon energies, this particular configuration indicates that thermal hard X-ray emission is observed hottest near the X-point, and progressively cooler with increasing distance from the X-point. This result was interpreted in terms of a current sheet formed above the flare loop location, as expected in the standard flare model. This observation can be considered as the first direct localization of a current sheet in a flare, and thus provides strong support for the standard model.

6. Footpoint motion and reconnection rate

The footpoint motion of hard X-ray sources or flare ribbons reveals the evolution and progression of coronal magnetic reconnection sites, because there is a direct magnetic mapping from the reconnecting X-points to the footpoints due to the low plasma β -parameter in the solar corona. It is therefore interesting to monitor the footpoint motion in order to obtain information on the location and rate of magnetic reconnection. In the standard 2D flare model it is expected that the X-point progressively rises to

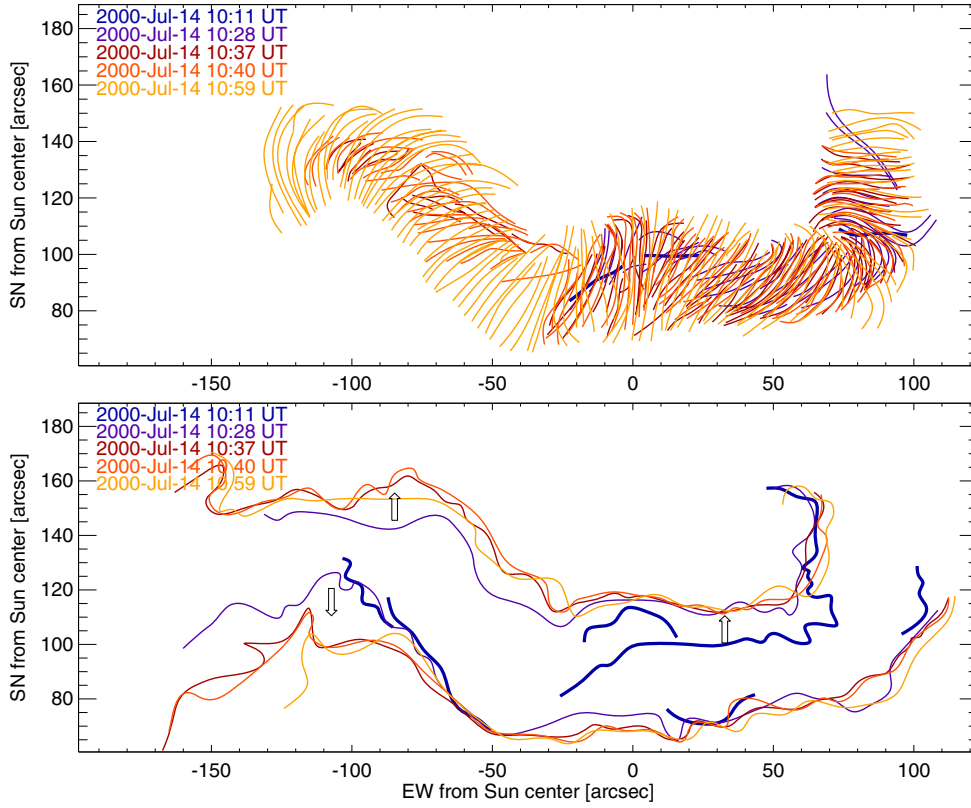


Figure 5. Top: Tracings of individual flare loops from TRACE 171 Å images of the Bastille-Day flare 2000 July 14, 10:11 UT, rendered with different colors during 5 subsequent time intervals. Note the evolution from highly sheared to less sheared loops. **Bottom:** The position of the two flare ribbons are traced from 171 Å images. Note the increasing footpoint separation with time (from Aschwanden 2004).

higher altitudes, and thus consequently the footpoint separation should increase with time. This is observable in some cases, for instance during the Bastille-Day 2000 flare (Fig. 5, bottom frame). However, because large flares always show double ribbons that can have a considerable length along the neutral line, up to 200,000 km (in the Bastille-Day 2000 flare), 3D flare models are required. Magnetic reconnection sometimes progresses along the neutral line (zipper effect), and thus the motion of footpoints and coronal hard X-ray sources is observed to move systematically parallel to the ribbons (e.g., Krucker *et al.* 2003). The motion parallel to the flare ribbons can be so strict that no perpendicular motion of footpoints is observed at all, which suggests that the reconnection progresses horizontally rather than vertically (Grigis & Benz 2005).

Based on reconnection models, the energy release rate can be written as the product of the Poynting flux S into the reconnection region and the area A of the reconnection region,

$$\frac{dE}{dt} = SA = 2 \frac{B_c^2}{4\pi} v_i A, \quad (3)$$

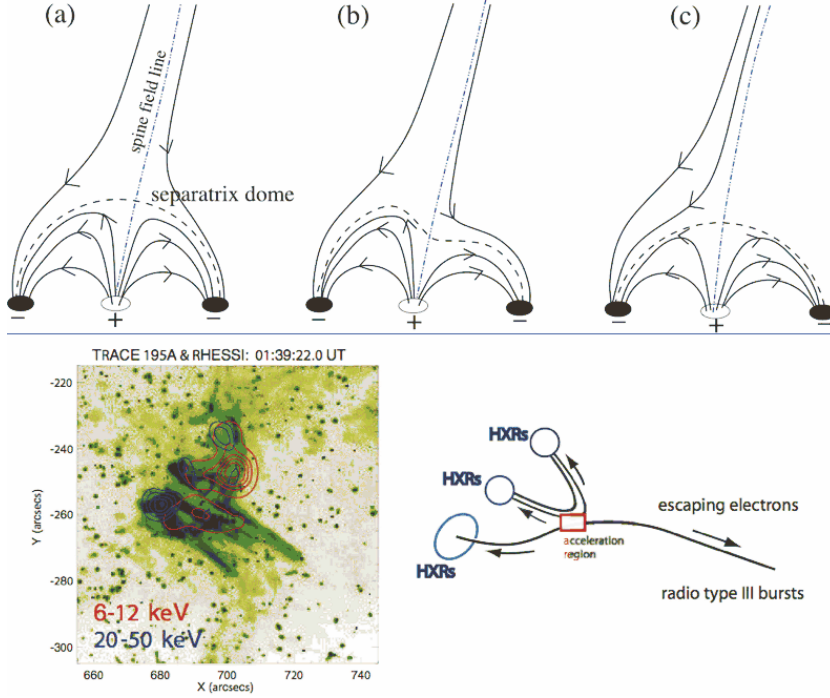


Figure 6. **Top:** Magnetic topology inferred for the 1993 May 3, 23:05 UT, flare, suggesting reconnection at a 3D nullpoint where the spine field line intersects the separatrix dome (Fletcher *et al.* 2001) **Bottom:** A similar topology is observed in a flare observed with RHESSI and TRACE (Krucker *et al.* 2004).

where B_c is the magnetic field strength in the corona and v_i is the inflow velocity into the reconnection region. Using the flare ribbon motion to estimate the lateral inflow velocity v_i and the extrapolated photospheric magnetic field to estimate the coronal value B_c , an approximate correlation was found between the estimated energy release rate and the observed hard X-ray or microwave fluxes (Asai *et al.* 2004).

7. Evidence for tripolar magnetic reconnection

There is a class of magnetic topologies that involve reconnection between an open field line and a closed field line, which in the simplest case corresponds to a “tripolar” configuration. After reconnection, one footpoint of the closed loop becomes the footpoint of the open field line. In more complex 3D topologies, an isolated polarity is surrounded by a region of opposite polarity, which creates a coronal nullpoint at the intersection of the separatrix dome and the vertical spine field line (Fig. 6, top). Flares that are consistent with such tripolar topologies (in a 2D plane) have been observed with TRACE (Fletcher *et al.* 2001) and with RHESSI (Krucker *et al.* 2004). A by-product of such tripolar flares is the appearance of linear jets that are detectable in soft X-rays and EUV, as well as escaping electron beams that produce radio type III bursts (Fig. 6, bottom).

8. Displaced electron and ion acceleration sources

The thick-target model assumes that nonthermal particles precipitate to the chromosphere and produce collisional bremsstrahlung there, so we expect that the hard X-ray footpoint sources are more or less cospatial at all high energies, except for a possible difference in chromospheric altitude. If ions are accelerated in the same source as the electrons, we expect also the gamma-ray emission resulting from nuclear de-excitation lines to be cospatial with the hard X-rays. It was therefore a big surprise that RHESSI observations for the first time revealed disparate footpoint sources for hard X-rays and 2.2 MeV gamma rays, which is produced by the capture of neutrons, produced by accelerated protons and heavier nuclei (C, N, O, Fe). A displacement of $\approx 20''$ was observed during the 2002 July 23, 00:30 UT, flare (Hurford *et al.* 2003). Similar displacements were also observed in other flares (Hurford *et al.* 2006). This came as a surprise and it was speculated that electrons and ions are either accelerated at different coronal sites (Emslie 2004), or that they follow different trajectories. No convincing explanation has been worked out yet.

9. Conclusions

TRACE and RHESSI continue to reveal new insights into solar flares. The following new results are briefly summarized in this paper: diagnostics on the multi-thermal flare plasma (sections 2, 3), the vertical structure of the chromosphere at flare loop footpoints (section 4), the complex magnetic topologies during flares that cannot be explained by the simple standard 2D model (section 5), the evolution and spatial progression of magnetic reconnection and the evolution of the reconnection rate (section 6), and separate localization of electron and ion acceleration sites and trajectories (section 8). We live in a golden era of vibrant progress in solar flare physics.

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X-ray Emission from Solar Flares

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Abstract. Solar X-ray Spectrometer (SOXS), the first space-borne solar astronomy experiment of India was designed to improve our current understanding of X-ray emission from the Sun in general and solar flares in particular. SOXS mission is composed of two solid state detectors, *viz.*, Si and CZT semiconductors capable of observing the full disk Sun in X-ray energy range of 4–56 keV. The X-ray spectra of solar flares obtained by the Si detector in the 4–25 keV range show evidence of Fe and Fe/Ni line emission and multi-thermal plasma. The evolution of the break energy point that separates the thermal and non-thermal processes reveals increase with increasing flare plasma temperature. Small scale flare activities observed by both the detectors are found to be suitable to heat the active region corona; however their location appears to be in the transition region.

Key words. Solar flares—X-ray detectors—X-ray line emission and continuum—break energy—microflares.

1. Introduction

The Solar X-ray Spectrometer (SOXS) mission (Jain *et al.* 2000a, 2000b, 2005, 2006a, b) was launched onboard an Indian geostationary satellite namely GSAT-2 on 8 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) payloads. 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 context to energy transport and acceleration time scales of particles. The energy resolution revealed by Si PIN detector is sub-keV. The SLD payload is designed and developed at the Physical Research Laboratory in collaboration with ISRO Satellite Centre (ISAC), Bangalore, and Space Application Centre (SAC), Ahmedabad.

The X-ray energy spectrum from a typical large solar flare is dominated by soft X-ray line and thermal (free-free) bremsstrahlung emission at $\varepsilon \approx 1\text{--}20\text{ keV}$, and collisional bremsstrahlung of non-thermal electrons at $\varepsilon \approx 20\text{--}1000\text{ keV}$ (Jain *et al.* 2000a, 2000b; 2005). The 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 of the SLD and sub-keV energy resolution of Si PIN detector allows measuring of the intensity and mean energy of the Fe and Fe/Ni line features at approximately 6.7 and 8 keV respectively (Phillips 2004; Jain *et al.* 2005, 2006b; Phillips *et al.* 2006) as a function of time in all classes of flares. The Fe-line complex is due mostly to the 1s-2p transitions in He-like and H-like iron, FeXXV and FeXXVI respectively, with associated satellite lines. The weaker Fe/Ni line feature at ~ 8 keV made up of emission from He-like nickel and more highly excited FeXXV ions is also evident in the more intense flares (Phillips 2004; Phillips *et al.* 2004).

The detailed investigation by Jain *et al.* (2006b) of X-ray spectra of M-class flares observed by Si detector of SLD showed that peak energy and equivalent width (w) of Fe line feature vary as a function of temperature of flare plasma. Further, they also showed that the whole energy spectrum of 4–20 keV cannot be fitted with single temperature approximation. They used two temperatures to obtain the best fit for line emission and continuum. The equivalent width of the Fe-line feature was observed to exceed the theoretical one in the temperature range of ≈ 20 –60 MK, which could indicate multi-thermal flare plasma or higher iron abundance. The ratio of the 6.7 keV Fe to the 8.0 keV Fe/Ni line fluxes is therefore expected to decrease progressively with higher plasma temperature (Phillips 2004).

The collisional bremsstrahlung produced by non-thermal electrons in many flares found to be very strong over thermal bremsstrahlung, and their contribution vary over-flare duration showing the break in the energy spectrum. However, in view of our currently poor knowledge of low energy cut-off of collisional bremsstrahlung we are not able to estimate precisely the number of electrons that participated in non-thermal X-ray emission. On the other hand, the poor spectral resolution of X-ray detectors put a constraint to measuring precisely the break energy point between thermal and non-thermal contributions. The improved energy resolution of sub-keV and 1.7 keV of Si and CZT detectors respectively in their dynamic energy range over the earlier detectors provide an opportunity to measure the break energy point more convincingly.

The heating of solar corona is a long enduring problem in solar physics and many investigators have proposed micro and nanoflares as potential candidates for heating the active region corona. Both Si and CZT detectors of SOXS mission have observed a large number of small scale flare activity during the last four years. The signatures of Fe-line feature emission in micro flares as evident by Si and CZT detectors support the concept of small scale flares as a source of coronal heating.

Therefore, the purpose of this paper is to present the potentialities of Si and CZT detectors of SOXS mission in context to quantifying X-ray emission characteristics of solar flares in general, and improving our current understanding of thermal and non-thermal mechanisms and the role of microflares in heating the active region solar corona in particular.

2. Observations

2.1 Instrumentation

The instrumentation of the SLD payload, its in-flight calibration and operation has been described by Jain *et al.* (2005, 2006a, 2006b). The SLD payload is functioning satisfactorily onboard the GSAT-2 spacecraft and so far more than 450 flares of importance greater 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 4–60 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 over the earlier detectors used for solar flare research in this energy range. However, the spectral resolution of CZT detector is 1.7 keV 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 operating temperature of both 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 03:55 and 06:30 UT everyday. However, after 06:30 UT, the temperature on the detectors exceeds the limit to cool down by thermoelectric cooler. Thus SLD payload provides uninterrupted data for ~ 2.5 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 6 June 2003.

2.2 Dataset

The temporal data, i.e., intensity (counts/s) as a function of time is revealed in four energy bands, 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. We selected flares of *GOES* importance class B, C and M for the current study of different objectives as described in the text.

2.3 Temporal and spectral evolution

In Fig. 1, we show the temporal evolution of 14 August 2004 and 27 July 2005 flares made from the spectral data of the Si and CZT detectors. The time resolution for spectral mode observations during quiet period is 3 s but during flare mode it is 100 ms, which however, is restricted to 287.5 s only in view of telemetry constraints. The intensity (counts/cm²/s/keV) of the light curves is shown in 4.1–10, 10–15 and 15–25 keV bands from Si detector, while in 25–35 and 35–55 keV bands from CZT detector. It may be noted that the 14 August 2004 flare is rising in the beginning as a slow heating of the plasma, and then impulsive transport followed by long enduring thermal component, while the 27 July 2005 flare is a gradual long enduring flare. The onset of flare is delayed in higher and higher energy channels but peaks earlier than previous bands, which suggests that radiative cooling is the dominant mechanism in cooling the flare plasma.

The spectral evolution of these two flares 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. It may be noted that the Fe and Fe/Ni lines are unambiguously visible in the count spectra at ~ 6.7 and ~ 8 keV respectively, however their intensity and width vary over time. The count spectra are de-convoluted over the instrumental response to obtain the photon spectra.

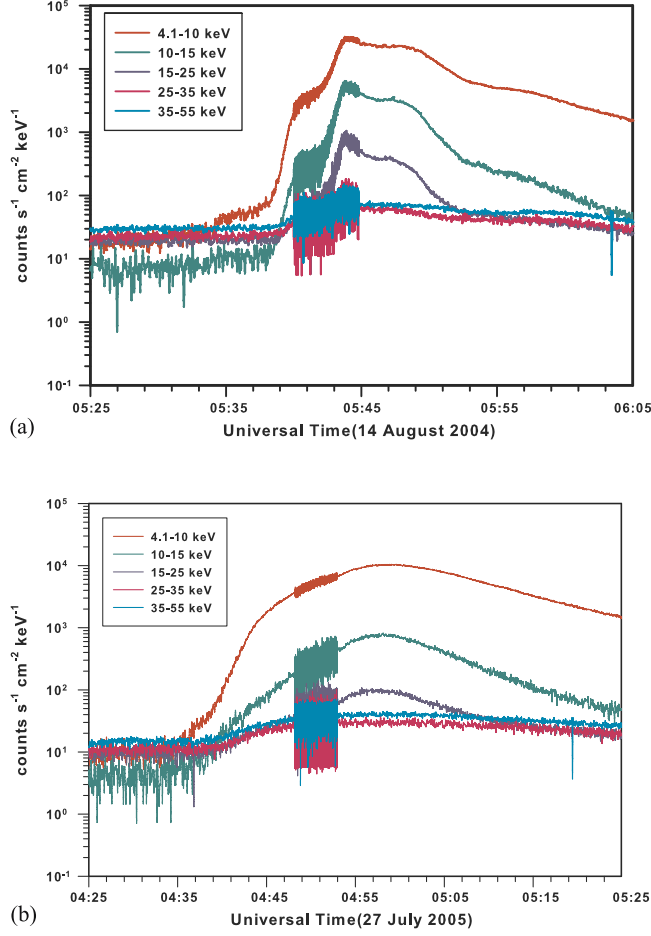


Figure 1(a & b). Light curves of the 14 August 2004 and 27 July 2005 solar flares as recorded by Si and CZT detectors of SLD/SOXS mission. 100 ms cadence during flare mode may be noted.

3. Analysis and results

We use the OSPEX (Object Spectral Executive) software package inside SolarSoft to analyse the data. The OSPEX is an object-oriented interface for X-ray spectral analysis of solar data. It is the next generation of SPEX (Spectral Executive) written by R. Schwartz in 1995. Through OSPEX, the user reads and displays the input data, selects and subtracts background, selects time intervals of interest, selects a combination of photon flux model components to describe the data, and fits those components to the spectrum in each time interval selected. During the fitting process, the response matrix is used to convert the photon model to the model counts to compare with the input count data. The resulting time-ordered fit parameters are stored and can be displayed and analyzed with OSPEX. The entire OSPEX session can be saved in the form of a script and the fit results stored in the form of a FITS file.

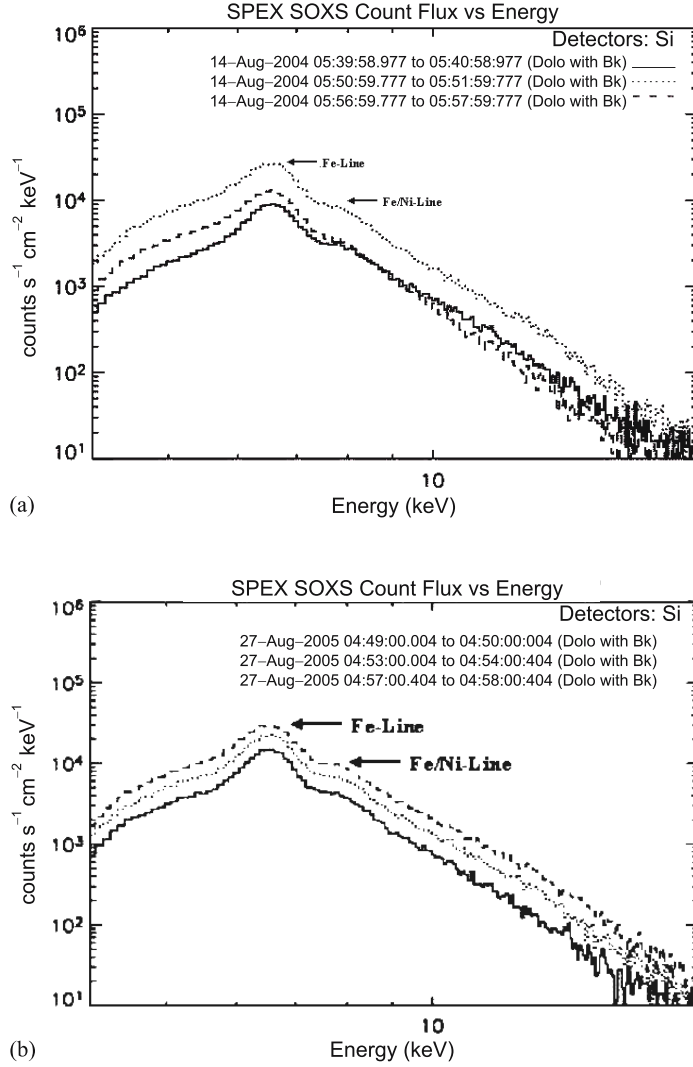


Figure 2(a & b). Spectral evolution of the 14 August 2004 and 27 July 2005 flares. The X-ray flares intensity (counts/cm²/s/keV) at three different intervals of time is shown as a function of energy. Fe and Fe/Ni line features are marked on the spectrum.

The OSPEX subroutines have been updated to undertake detailed temporal and spectral data analysis of Si and CZT detectors of SLD/SOXS mission. The raw data for spectral mode observations was first corrected for any spurious or false flare as well as for pre-flare background. The spectrum at a given time is made by integrating the spectra over an interval of 60 s period. The OSPEX enables us to look at the temporal evolution, spectrogram and spectral evolution. It also enables us to fit energy spectra using CHIANTI codes (Dere *et al.* 1997) for flare plasma diagnostics with the application of various thermal, line emission, multi-thermal, and non-thermal functions.

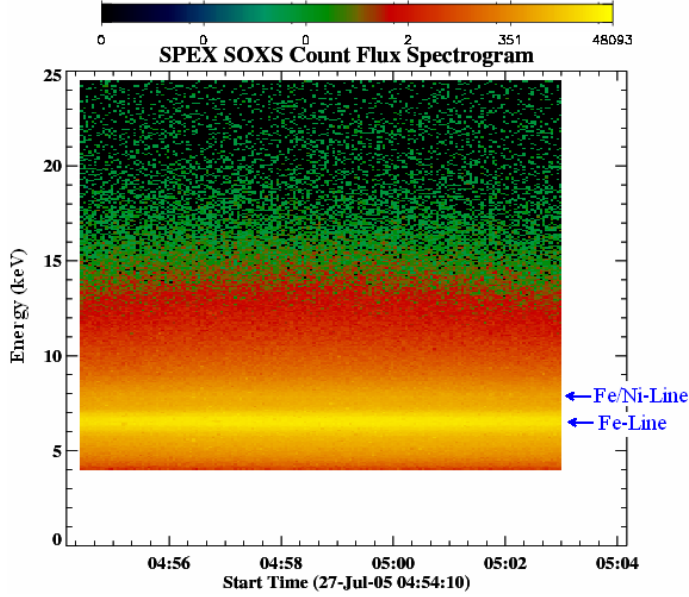


Figure 3. The spectrogram showing variation of Fe and Fe/Ni line features as a function of time during the flare of 27 July 2005.

3.1 X-ray emission from Fe line and Fe/Ni lines

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). It may be noted from Fig. 2 that the peak intensity, peak energy and area under the curve of the lines vary over time. In fact the plasma temperature and hence emission measure vary over time and these factors mainly control the shape of the line. The non-thermal contribution also plays a role and therefore needs to be considered as an important parameter including temperature and emission measure.

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, and similarly the Fe/Ni line feature in the range 7.5–9.2 keV (Phillips 2004, Jain *et al.* 2006b). It may be noted from the flare spectrogram shown in Fig. 3 that the Fe and Fe/Ni features including the peak energy and intensity vary over the flare evolution, which suggests that the abundance and centroid/peak energy of the emission line vary as a function of temperature.

We analyzed 32 and 45 spectra of 14 August 2004 and 27 July 2005 flares respectively. We use forward fitting program in OSPEX to fit the count spectra where instrumental response matrix for Si and CZT detectors is incorporated. The important aspect of our spectra-fit analysis is consideration of multi-thermal power-law function, which assumes differential emission measure (DEM), in addition to line emission function, and non-thermal power-law function in order to estimate the contribution of non-thermal electrons in the flare plasma. The DEM in the multi-thermal power-law is of the following form:

$$\text{DEM}(T) = a_0 \cdot \left(\frac{2}{T}\right)^{\gamma_{MT}}, \quad (1)$$

where a_0 is differential emission measure at temperature $T = 2 \text{ keV}$ in units of $10^{49} \text{ cm}^{-3} \text{ keV}^{-1}$ and γ_{MT} is power-law index for calculating the differential emission measure at the temperature T obtained from multi-thermal power law fit of the observed spectrum.

The multi-thermal power-law function enables us to measure the emission measure, the minimum (assumed 5.8 MK) and maximum temperature of the flare plasma, power-law index and the abundance of Fe-line relative to corona. The line emission Gaussian function provides integrated intensity, centroid energy of the line and line width. On the other hand, the non-thermal function allows us to measure power-law index below and above break energy. In Fig. 4, we show a typical example of count spectra-fit with the application of the above functions.

It may be noted from Fig. 4 that non-thermal bremsstrahlung also contributes considerably to the evolution of Fe and Fe/Ni line features in addition to significant

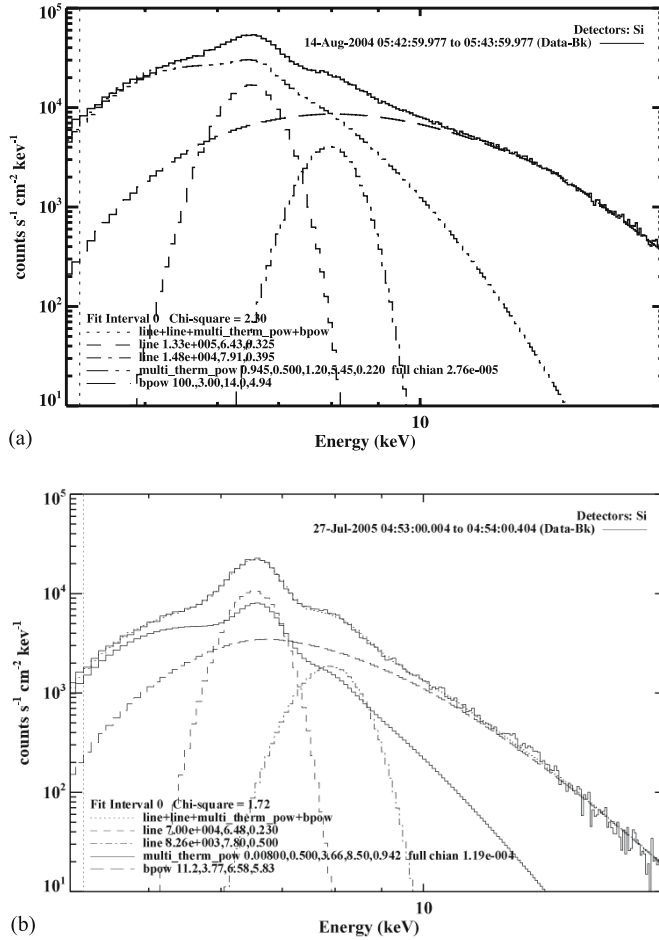


Figure 4. (Continued)

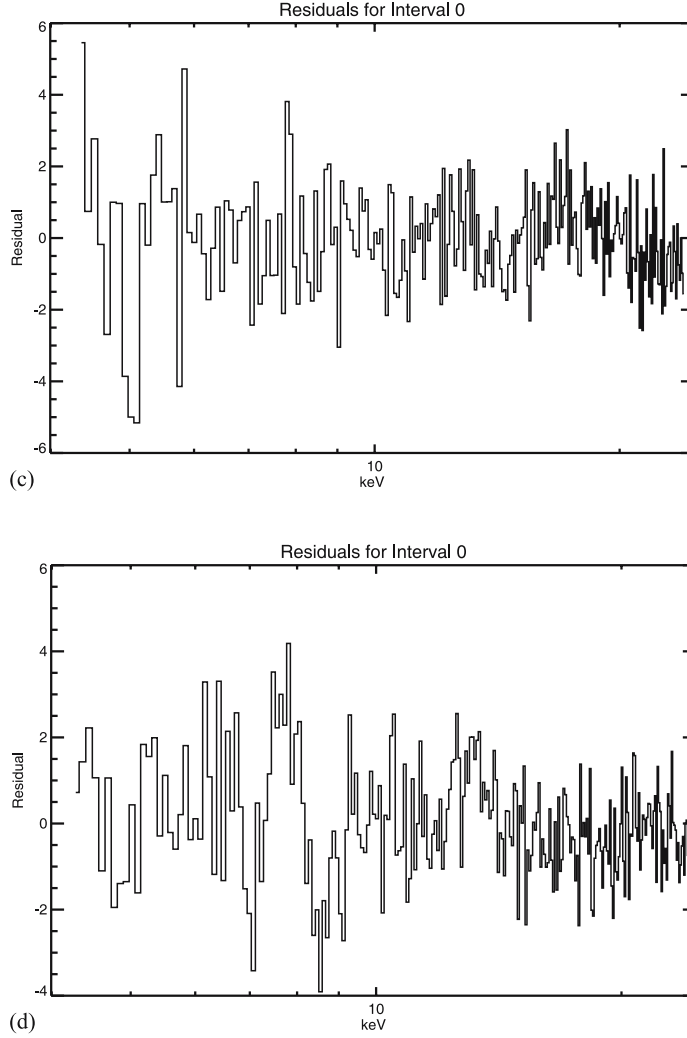


Figure 4(a–d). Fit count spectra (a & b) observed near peak time of 14 August 2004 and 27 July 2005 flares considering line Gaussian, multi-thermal and broken power-law functions using Chianti codes. Residuals as a function of energy for the fit spectra are also shown in (c) and (d).

contribution by thermal processes inside the flare plasma. Once the count spectrum is fitted, it can be de-convoluted over the instrumental response to obtain the photon spectrum which is generated using the fitted model count spectrum. We do not use the pulse piled up spectra for analysis to avoid false temperature results. To obtain a reasonably satisfactory model fit to the count spectra we fit the parameters taking in consideration the following conditions:

- Parameters should give satisfactory values for the reduced chi-squared ($\chi^2 < 3$).
- The temperature and emission measure should have a consistent time history, i.e., their variation should not be very large from one to the next spectra.

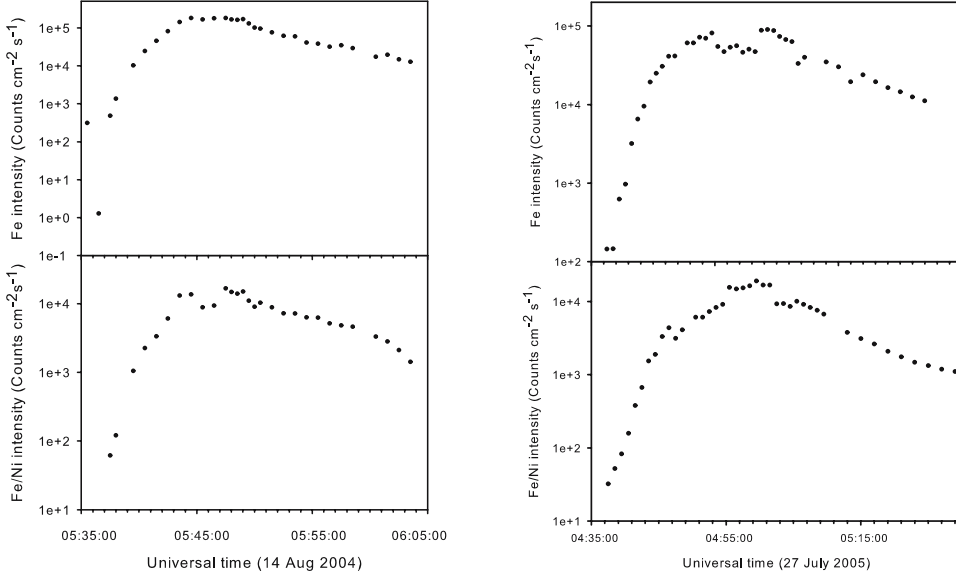


Figure 5. The intensity of Fe and Fe/Ni line features varying over time in 14 August 2004 and 27 July 2005 flares.

In Fig. 5, we show the integrated line intensity variation of Fe and Fe/Ni line features as a function of time measured in 14 August 2004 and 27 July 2005 flares. It may be noted that the line intensity follows by and large the flux intensity of the flare in 4–10 keV. However, line intensity of Fe and Fe/Ni features is found to increase as a function of temperature measured in these two flares as shown in Fig. 6. It may be noted that the increase in the intensity of the line feature with temperature is exponential up to 25 MK and then it remains to almost the same level until ~ 80 MK. Based on the analysis of more than 10 flares, Jain *et al.* (2006a) showed that equivalent width (abundance) of Fe line feature increases exponentially up to 25 MK and remains stable up to 45 MK. However, it reduces with temperature exceeding 45 MK. Perhaps, a similar result may be observed if analysis were conducted for large number of flares.

3.2 X-ray emission from flare plasma

In Fig. 7, we show the variation of flux in 10–20 keV, temperature and emission measure over time for these two flares of 14 August 2004 and 27 July 2005. It may be noted when temperature of the flaring plasma is decreasing the emission measure is increasing.

If the flare is dominated by thermal emission mechanism then evolution of flux follows the evolution of temperature as may be noted in the case of 27 July 2005 flare (cf. Fig. 7 – right side). On the contrary, at the onset of the 14 August 2004 flare, temperature is not high rather DEM is very high and remains between 10^{47} and $10^{50} \text{ cm}^3 \cdot \text{keV}^{-1}$ during rise phase. This suggests that non-thermal bremsstrahlung was the principal mechanism for onset of the flare, which perhaps was generated by acceleration of particles causing the large ionization, and hence producing very high electron density revealed by large DEM. We estimated volume of the flare from

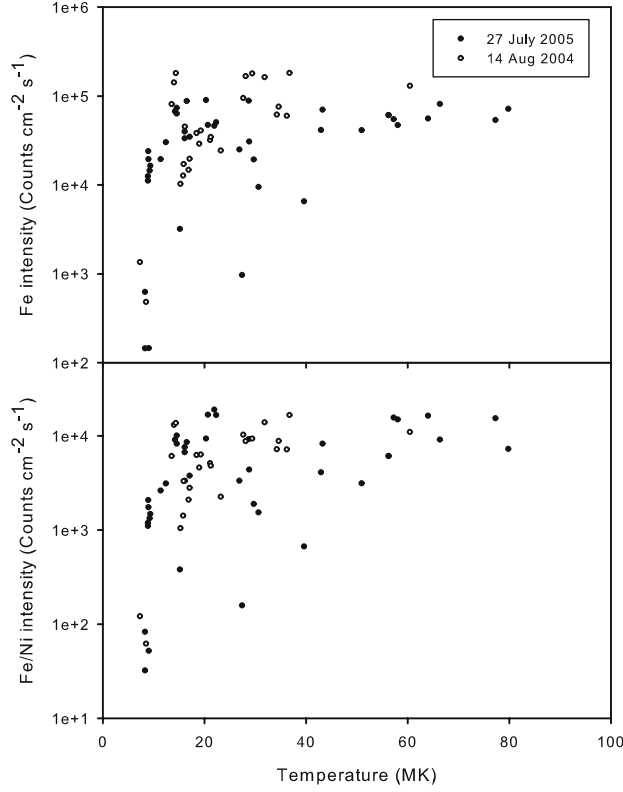
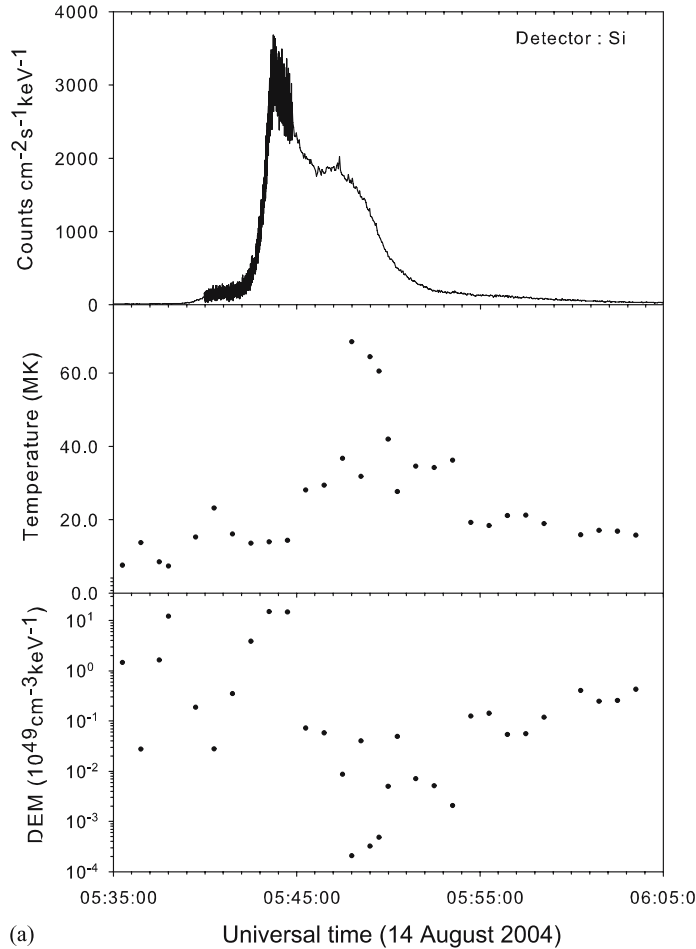


Figure 6. Integrated line intensity of Fe and Fe/Ni features as a function of temperature measured in 14 August 2004 and 27 July 2005 flares.

$H\alpha$ images during rise phase varying between 10^{27} and 10^{28} cm^3 suggesting electron density (N_e) to be $> 10^{11} \text{ cm}^{-3}$. After the peak of the flare around 05:44 UT the temperature of the flare begins to grow to reach more than 70 MK in the superhot phase of the flare around 05:47 UT. The temperature falls in the decay phase. On the other hand, temperature profile of 27 July 2005 flare appears similar to count flux. However, DEM falls exponentially from 10^{49} to $10^{42} \text{ cm}^3 \cdot \text{keV}^{-1}$ during rise phase of the flare and then increasing during decay phase suggesting that most of the X-ray emission is caused by heating of the plasma. The measurement of temperatures, emission measure at 2 keV, and multi-thermal power-law index by spectra fit enabled us to justify the relation (1) that DEM decreases as power-law with increase in temperature as shown in Fig. 8.

The negative power-law index below (γ_1) and above (γ_2) break energy measured from broken power-law fit of the spectrum in simultaneous to multi-thermal and line emission is shown in Fig. 9 as a function of plasma temperature (T_e). The power-law index below break energy represents thermal part of the flare spectrum and therefore it increases with temperature as may be noted from Fig. 9. On the contrary, power-law index above (γ_2) break energy representing non-thermal part of the spectrum is very high at lower temperature and decreases with T . In the case of the 14 August 2004 flare, the γ_2 was higher during rise phase in context to non-thermal onset of the flare, while it was very low in the case of 27 July 2005 flare as the onset was thermal.

**Figure 7.** (Continued)

3.3 Break energy

The determination of the low-energy cut-off to nonthermal electron distributions is critical to the calculation of the nonthermal energy in solar flares. The most direct evidence for low-energy cutoffs is flattening of the power-law, nonthermal X-ray spectra at low energies. However, because of the plasma preheating often seen in flares, the thermal emissions at low energies may hide such spectral flattening of the nonthermal component. We select the 14 August 2004 and 27 July 2005 flares to estimate the break energy (cut-off energy between thermal and non-thermal emission) considering first the concept of “early impulsive flares”, in which the ≥ 25 keV hard X-ray (HXR) flux increase is delayed by less than 30 s after the flux increase at lower energies. Thus, the plasma preheating in these flares is minimal, so the nonthermal spectrum can be determined to lower energies than in flares with significant preheating. As shown earlier in section 3.2, the flare of 14 August 2004 is impulsive in the early stage, while that of 27 July 2005 began with thermal emission. In these events, the break energy of the multi-thermal power-law and non-thermal power-law fit to the HXR spectra lies

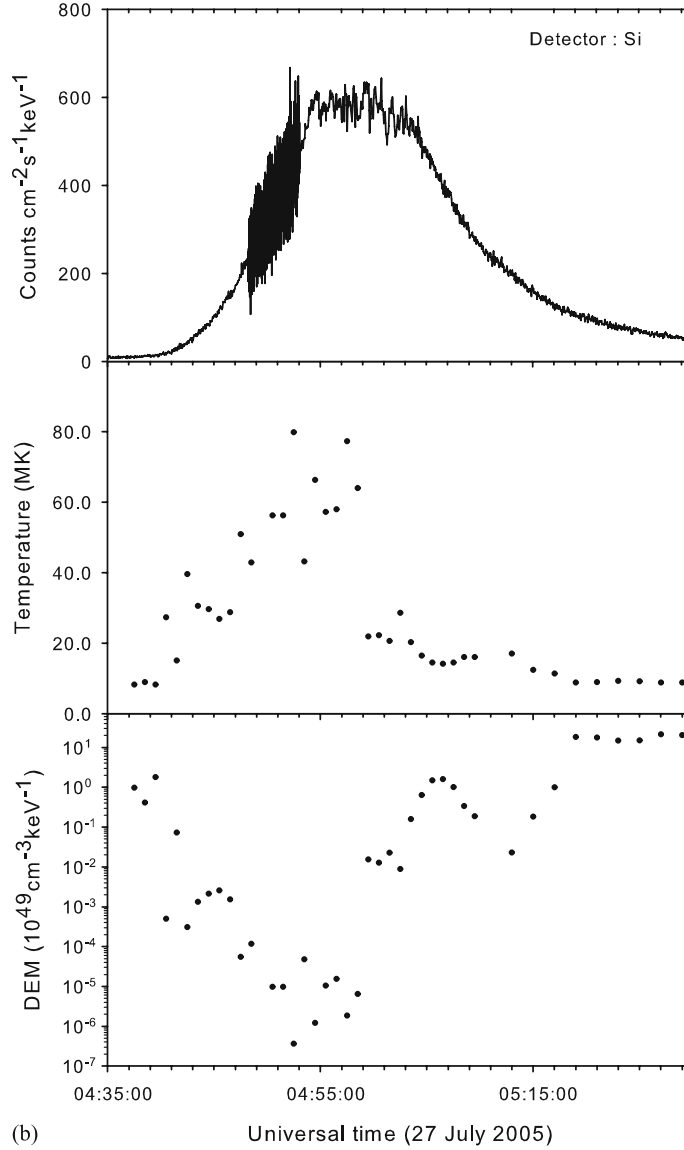


Figure 7(a & b). Variation of count flux in 10–20 keV (top panel), temperature (middle panel) and differential emission measure (bottom panel) as a function of time in the 14 August 2004 and 27 July 2005 flares.

in the range 5–20 keV, significantly lower than the value we have seen for other flares that do not show such early impulsive emissions. We show in Fig. 10 the evolution of break energy in the 14 August 2004 and 27 July 2005 flares measured from the spectra observed by Si detector with ± 0.8 keV energy resolution. It may be noted that the break energy in the 14 August 2004 flare is very low in the beginning of the flare due to non-thermal onset of the flare, and it slowly increases with time as the temperature of the flare also increases. While in the case of the 27 July 2005 flare event it is high in

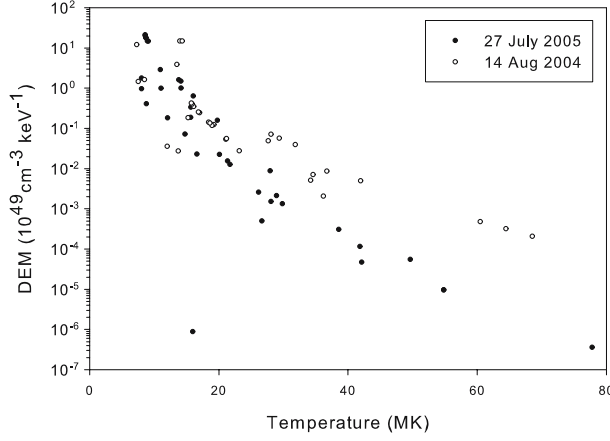


Figure 8. The differential emission measure (DEM) as a function of temperature measured in the 14 August 2004 and 27 July 2005 flares observed by Si detector of SOXS mission.

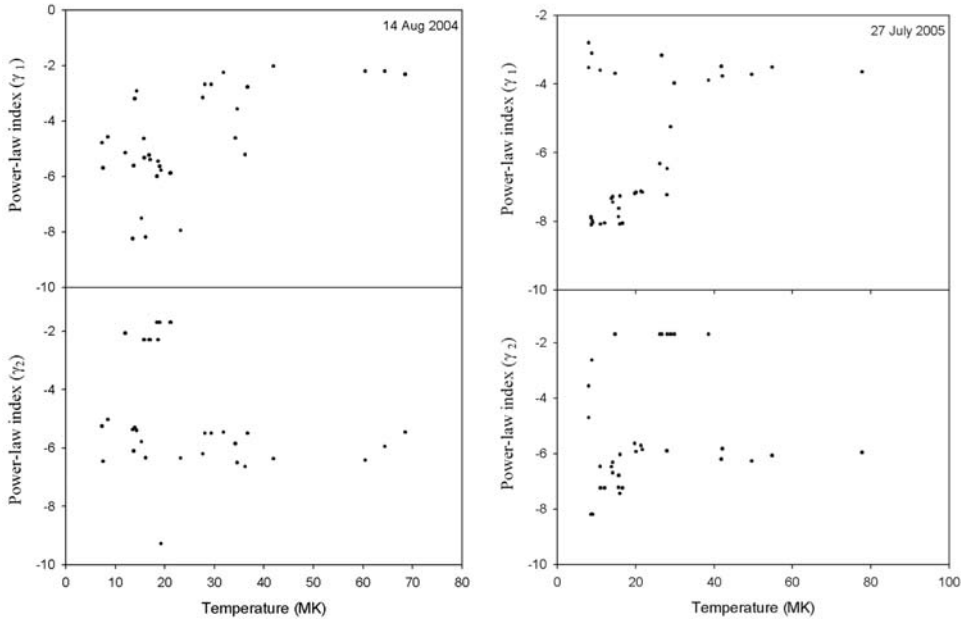


Figure 9. Variation of broken power-law index below (γ_1 – top panel) and above (γ_2 – bottom panel) break energy as a function of flare plasma temperature measured in the 14 August 2004 and 27 July 2005 flares.

the early phase due to thermal onset of the flare, however, it decreases with time as temperature goes down but electron density had already increased as revealed from high differential emission measure.

In Fig. 11 variation of break energy as a function of temperature of the flare plasma is presented. It may be noted that the break energy increases with temperature up to 40 MK. We do not have enough points above 40 MK to suggest that the break energy

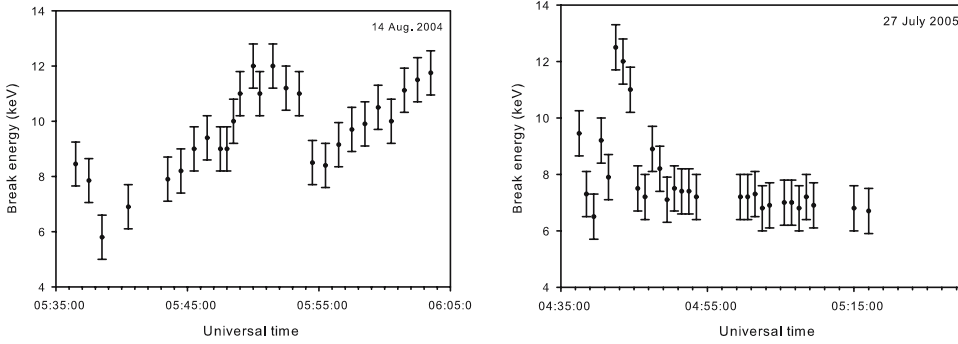


Figure 10. Break energy as a function of time measured by Si detector with 0.8 keV FWHM in the 14 August 2004 and 27 July 2005 flares.

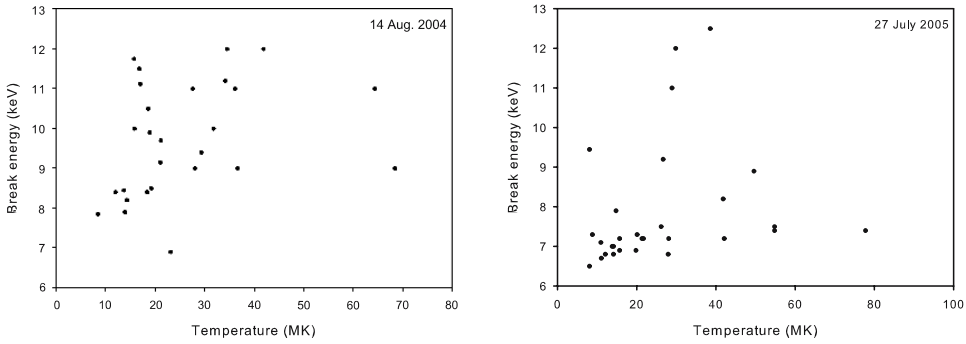


Figure 11. Break energy as a function of flare plasma temperature in the 14 August 2004 and 27 July 2005 flares. Estimation of break energy is restricted to ± 0.8 keV, resolution of the Si detector.

falls above this temperature though it appears in both flares. The increase of break energy with temperature suggests that low energy cut-off of non-thermal mechanism shifts to higher energy due to increasing contribution of thermal bremsstrahlung at higher and higher temperatures (super-hot) of the flare plasma.

3.4 Microflares

Solar flares propel energetic particles into space and are thought to be the main source of heat pumping the Sun's outer atmosphere to about a few million degrees Celsius. But with the investigation of small scale flares called micro and nanoflares by YOHKO, SOHO and RHESSI it has been suggested that these tiny numerous microflares may influence solar atmosphere more than the solar flares. Now observations show that microflares a million times smaller are more frequent and may provide a major portion of the heat in the corona. The microflares are identified by their hard X-ray emission. However, there are observations of microflares being mainly thermal events, heating the Sun but not accelerating particles like larger flares (Jain *et al.* 2006c). If that were the case, they would produce more low-energy soft X-rays than high-energy hard

X-rays. Interestingly, a subset of microflares is responsible for type III radio bursts, which, however, do very little heating of the solar corona. Instead, the stream of high-speed particles they produce seems to jet unchecked out of the Sun at speeds up to one-third the speed of light, exciting radio oscillations at lower and lower frequencies as the particles pass through lower and lower density plasma. This probably has to do with the magnetic field in the region around the microflare, since particles are tied to the field lines and have to run along them, we think that for normal microflares, the particle acceleration occurs in a closed magnetic region so the electrons cannot get away; they do more heating that way. In type III bursts, the electrons are accelerated in an open magnetic field, and they have an easy way to escape, so they do some less of heating. Mirzoeva (2006) using the data from Interball–Geotail project for the period September through December 1995 derived the intensity distribution of microflares and found correlations between the daily mean peak fluxes of X-ray bursts from microflares of various classes and the daily mean values of the thermal background of the solar corona. He suggested that microflares play a significant role in heating the solar corona. Jain *et al.* (2006a) showed that the minimum temperature required for the formation of Fe and Fe/Ni line features are 8 and 14 MK respectively. In view of high Z of these elements they settle down very fast against gravitation of the Sun as soon as the temperature of the corona goes below these temperatures.

We investigate low-intensity microflares in the soft and medium hard component (4–20 keV) of the solar X-ray radiation and explore the possibility of the presence of Fe and Fe/Ni line features to probe their role in heating the corona.

Shown in Fig. 12 are the spectrograms of the microflares observed on 1 May 2004, 27 April 2007 and 3 June 2007 by Si detector, which unambiguously represent the presence of Fe line feature. In all three microflares though weak, signatures of Fe/Ni line feature are also visible further indicating that temperatures in these microflares could reach to 14 MK. Presence of these line features in microflares show their strength to heat the active region corona to 8 MK or more. On the other hand their spectra fit provide an evidence of the presence of thermal plus non-thermal components indicating heating of the plasma as well as the particle acceleration inside the magnetic loops.

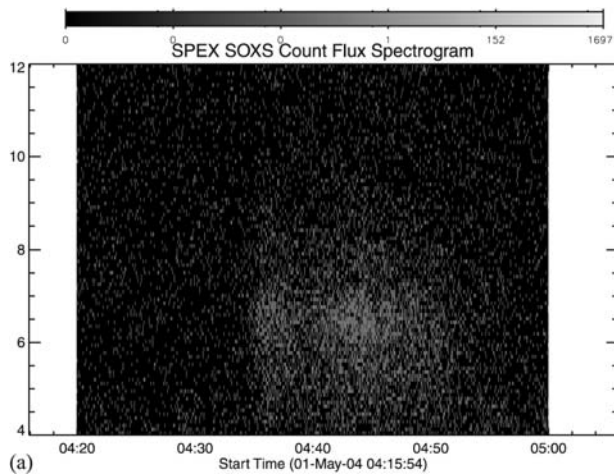


Figure 12. (Continued)

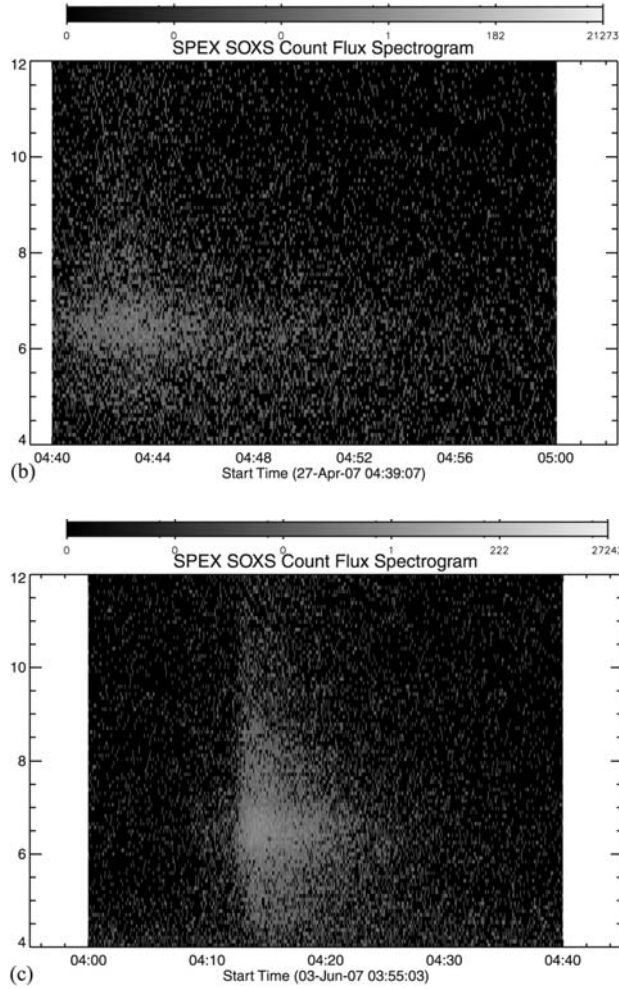


Figure 12(a-c). X-ray spectrograms of microflares observed by Si detector on 1 May 2004, 27 April and 3 June 2007. The intense Fe line feature around 6.7 keV during the flare interval may be noted in all flares, while faint signature of Fe/Ni line feature around 8 keV may also be seen in 27 April and 3 June 2007 microflares.

In Fig. 13, we show the spectra integrated over one minute during rise phase of the microflares, viz., 27 April 2007 and 3 June 2007. It may be noted that spectra of 27 April 2007 is best fitted by multi-thermal while that of 3 June 2007 by multi-thermal plus non-thermal assumptions. However, sometimes in small microflares ($< B5$) we are not able to see the signature of Fe line feature perhaps due to the sensitivity of the detector limiting to this intensity. In future if detectors with higher sensitivity are flown then not only micro but also nanoflares may show the presence of Fe line feature, which may further emphasize the possibility of these small scale activities to be responsible for heating the solar corona.

The count flux, temperature and differential emission measure as a function of time of these three microflares under current investigation are shown in Fig. 14. It may be

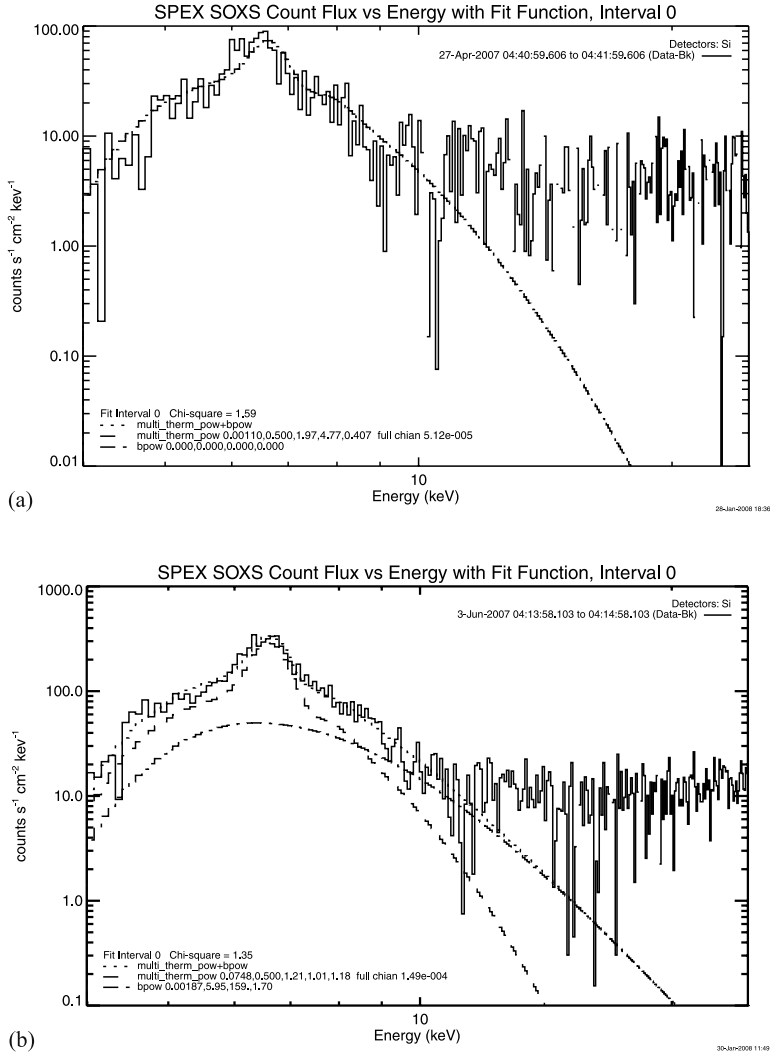


Figure 13(a & b). Count spectra fit by multi-thermal and broken power-law assumption of the 27 April and 3 June 2007 microflares during their rise phase.

noted that the microflare of 1 May 2004 is weak and slow rising while the 27 April 2007 is medium intense, and 3 June 2007 appears to be a fast rising microflare and stronger than the other two flares. Power-law inverse relationship between temperature and differential emission measure may also be observed in the microflares.

A small pre-heating was observed in the 1 May 2004 microflare during which maximum temperature was around 18 MK. The temperature decayed until 04:40 UT and then started to increase with main microflare to reach maximum about 24 MK. However, it may be noted that emission measure was very small in this microflare ranging between 10^{38} and $10^{43} \text{ cm}^{-3} \cdot \text{keV}^{-1}$. This suggests that neither particle acceleration nor the ionization of the plasma was the case in this microflare, rather simple heating of the ambient plasma by increased temperature appears to be more appropriate. On

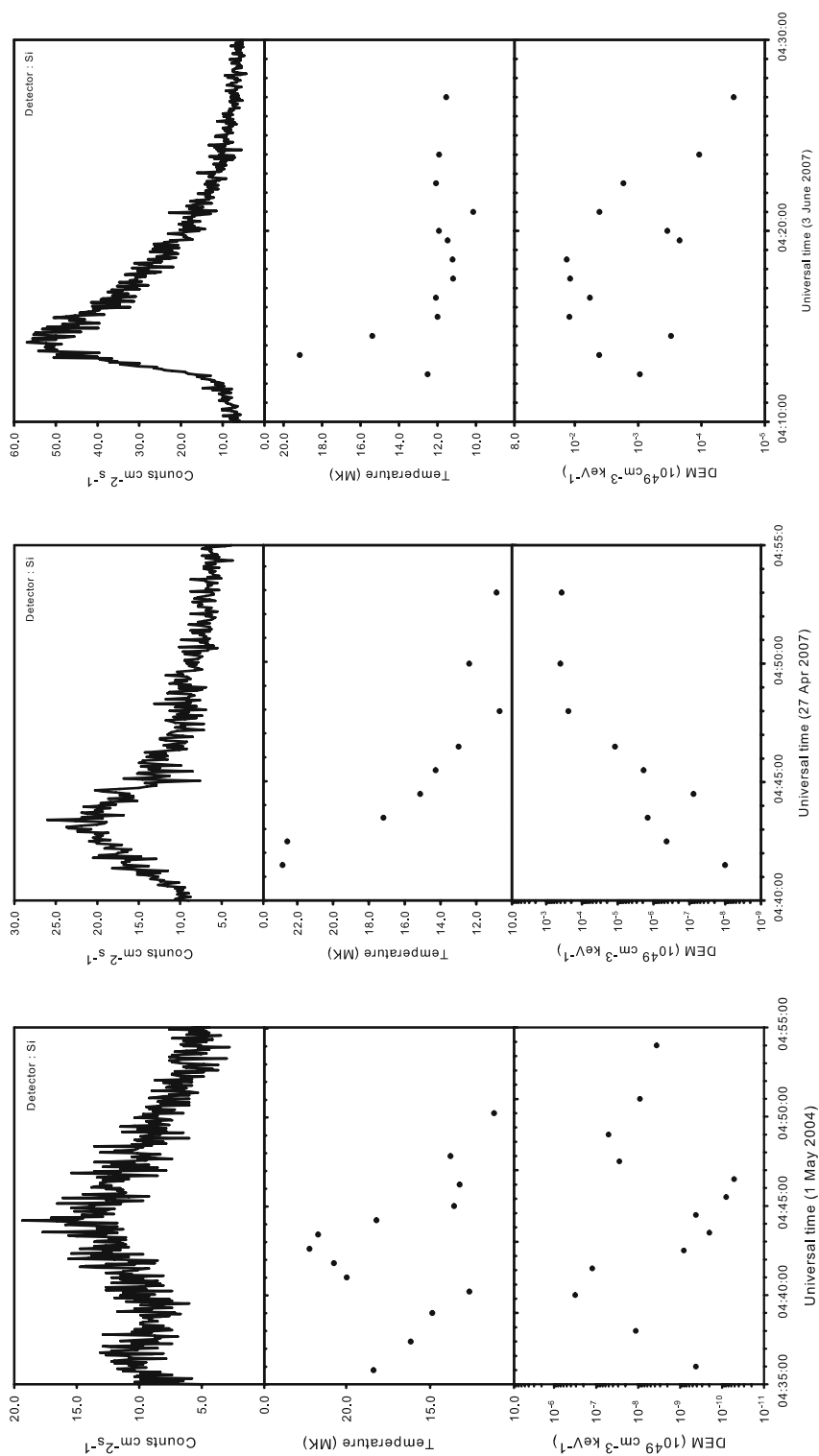


Figure 14. Variation of X-ray count flux, temperature and emission measure of flare plasma as a function of time in 1 May 2004, 27 April 2007 and 3 June 2007 microflares.

the other hand the case of 27 April 2007 microflare ionization by heating or particle acceleration in the closed loop structure appears as a suitable case in context to power-law increasing of emission measure from 10^{41} to $10^{46} \text{ cm}^{-3} \cdot \text{keV}^{-1}$ with decrease of temperature. The case of 3 June 2007, however, appears to us of hybrid configuration where plasma heating by both temperature and particle acceleration was in the progress as may be noted from the thermal and non-thermal fit of the spectra shown in Fig. 14(b). The emission measure was observed to range between 10^{44} and $10^{47} \text{ cm}^{-3} \cdot \text{keV}^{-1}$, significantly higher than the other two microflares.

No signature of type III radio bursts were observed in association to these flares. Thus it appears that the magnetic reconnection was in closed loop structure in the lower corona or transition region causing heat by thermal as well as by accelerated particles, which thermalized during their gyration along the loops. However, recently Aschwanden (2007) showed that the energy released by microflares/nanoflares in the corona is not enough to heat the corona. He suggested that if such small scale flares are considered to heating upper chromosphere and/or transition region then they may also play a significant role in coronal heating.

4. Discussions and conclusions

In the current analysis of SOXS flare data, the spectra in the dynamic energy range 4–23 keV during flare interval were analyzed considering line Gaussian, multi-thermal power-law and broken power-law functions to fit the Fe-line and Fe/Ni-line features with Gaussian profiles at 6.7 keV, and ~ 8 keV respectively, and thermal and non-thermal spectra consisting of a continuum (free-free plus free-bound emission). The data analysis was performed using OSPEX/SolarSoft package that accounts for all known instrumental effects and allows for background subtraction. The multi-thermal power-law was used in view of varying emission measure as a function of temperature, in addition to line Gaussian and broken power-law to fit the spectra. The fits to the model spectra are most reliable for Si detector's spectra taken during the flares (as measured by the reduced χ^2).

Phillips *et al.* (2006) and Jain *et al.* (2006b) showed that the observed equivalent widths (w) of Fe line feature increase with T_e more sharply than is predicted by the theoretical curve, with a clear displacement to higher temperatures being indicated. These results are not so easily explained. If the isothermal approximation is a good one for spectra during these flares, as is indicated by the reduced χ^2 values, then possibly this behaviour could indicate the need for a correction to the ionization fractions of Mazzotta *et al.* (1998). This suggests to differential emission measure indicating to multi-temperature model instead of single temperature approximation that was considered so far in the X-ray spectral analysis of the flares. In this investigation intensity of the Fe and Fe/Ni line features was found to increase with temperature up to 45 MK. However, beyond 45 MK increase or decrease was not clear due to limited observations. Thus we propose to analyze more flares of similar intensity and by the same techniques to investigate behaviour of line intensity above 45 MK.

We found that 14 August 2004 flare started as non-thermal mechanism while 27 July 2005 as plasma heating process. However, the differential emission measure shows power-law decrease as a function of the flare plasma temperature (T_e). This was confirmed by break energy, which increases with temperature, as well as by the variation of spectral index (γ_1) below breaks energy, which is representative of free-free

continuum thermal bremsstrahlung emission, and increase with temperature. On the other hand the spectral index γ_2 , above break energy, representative of free-bound continuum, i.e., non-thermal emission reduces with temperature.

We found the signatures of Fe and Fe/Ni line features in microflares of varied intensity indicating flare plasma temperatures exceeding 14 MK required to form Fe and Fe/Ni line features. Their spectra analysis unambiguously showed temperatures more than 20 MK, and particle acceleration, however, apparently in closed loop configuration. Based on the evidence of line features and spectra fit by multi-thermal and broken power-laws we conjecture hybrid model for microflares with thermal plus non-thermal mechanisms. Incidentally, the energy released by such microflares is not found sufficient to heat the active region corona (Aschwanden 2007). However, the absence of type III radio bursts in association with the microflares studied by us indicate that they had very small and closed loops, and therefore originated either in upper chromosphere or in transition region where the “tectonic” photospheric random motion driven by subphotospheric convection is likely to cause small-scale reconnection processes in the transition region (Priest *et al.* 2002), which heat up the plasma at the footpoints of coronal loops and fill them up subsequently. Recent hydrodynamic simulations by Warren *et al.* (2002, 2003) employ an intermittent heating process localized at the loop footpoints, which produces a sequence of repeated heating pulses that are able to mimic the observed temperature evolution of coronal loops. As soon as the over-pressure starts to fill up the loop with hot plasma, the loop is expected to brighten rapidly in the soft X-rays first (during a filling-time interval that is in the order of a shock loop-crossing time, i.e., within minutes for typical coronal loops), without any detectable signal in the cooler EUV channels. Consequently, we would observe only the cooling phase, when the loop cools down from the soft X-rays through EUV. Such loop cooling phases from *Yohkoh* soft X-ray filters down to *TRACE* EUV filters have indeed been observed in detail for 11 cases, over time intervals of several hours, without any noticeable EUV signature of an initial heating phase (Winebarger & Warren 2005; Ugarte-Urra *et al.* 2006). The footpoint of a loop has been monitored in EUV during the initial rise in soft X-rays in detail in a single case (Fig. 6 in Ugarte-Urra *et al.* 2006), but the full loop was not detected in EUV ($T = 1.0\text{--}1.5$ MK) until it cooled down from the hotter ($T = 3.8$ MK) soft X-ray temperatures. Caveats of the “chromospheric evaporation scenario” for heating of quiescent loops are:

- the non-detection of the initial heating agent (small-scale reconnection events in the transition region) and
- the lack of observed blue-shifted upflows, which are probably to blame on insufficient time cadence, spatial, and spectral resolution.

We propose that plasma heating resulting from nanoflares and microflare events occurs in the transition region, rather than throughout the corona in higher altitudes. Recent imaging observations of EUV nanoflares and soft X-ray transient brightenings invariably show small-scale loops with length scales of $L = 2000\text{--}3000$ km that barely stick out of the transition region (Aschwanden & Parnell 2002). In the model of Parker (1988), where coronal field lines are more or less uniformly twisted along their lengths, heated plasma would spread along the coronal field lines and would manifest itself in long (probably unresolved) thin threads, for which we have no observational evidence. Thus we see strong support for the working hypothesis that the major heating

source responsible for a hot corona is located in the transition region rather than in the corona.

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X-ray Studies of Flaring Plasma

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Abstract. We present some methods of X-ray data analysis employed in our laboratory for deducing the physical parameters of flaring plasma. For example, we have used a flare well observed with Polish instrument RESIK aboard Russian *CORONAS-F* satellite. Based on a careful instrument calibration, the absolute fluxes in a number of individual spectral lines have been obtained. The analysis of these lines allows us to follow the evolution of important thermodynamic parameters characterizing the emitting plasma throughout this flare evolution.

Key words. Sun—flares—X-rays—spectroscopic plasma diagnostic.

1. Introduction

Often used and a simple method of coronal plasma parameters determination (such as mean temperature and density) is the lines/filters ratio technique. Although an isothermal assumption on which this method relies on is not fully applicable for many solar features, it is frequently used as a first approximation. In the present study, we incorporate also a more advanced analysis approach for a selected flare on 2002 October 4 (max. at 05:38 UT) well observed by Polish spectrometer RESIK. The nominal wavelength coverage of RESIK is 3.35 Å–6.1 Å in four independent channels. This range contains many emitting lines of Si, S, Ar, and K suitable for diagnostics of emitting plasma. A detailed description of the instrument, its operation and calibration is presented in the paper by Sylwester *et al.* (2005). Our selected event occurred on the solar disc (S18W08) and was of M4.0 class in *GOES* classification (1N in H α). The complementary data available for this flare are the 3 s *GOES* fluxes and the hard X-ray data from *RHESSI*. The *GOES* data have been used to determine the temperature and emission measure in the isothermal approximation. We have reconstructed *RHESSI* images at several times and for different energy bands using the CLEAN algorithm from the SolarSoft package. We assumed a constant volume of emitting region based on the contour profile drawn at 50% of maximum on *RHESSI* image (3–6 keV).

2. The data analysis

In Fig. 1 we present the four RESIK spectra observed during 50 s of the rise phase of the analysed flare. The identification of basic lines is given in the papers by

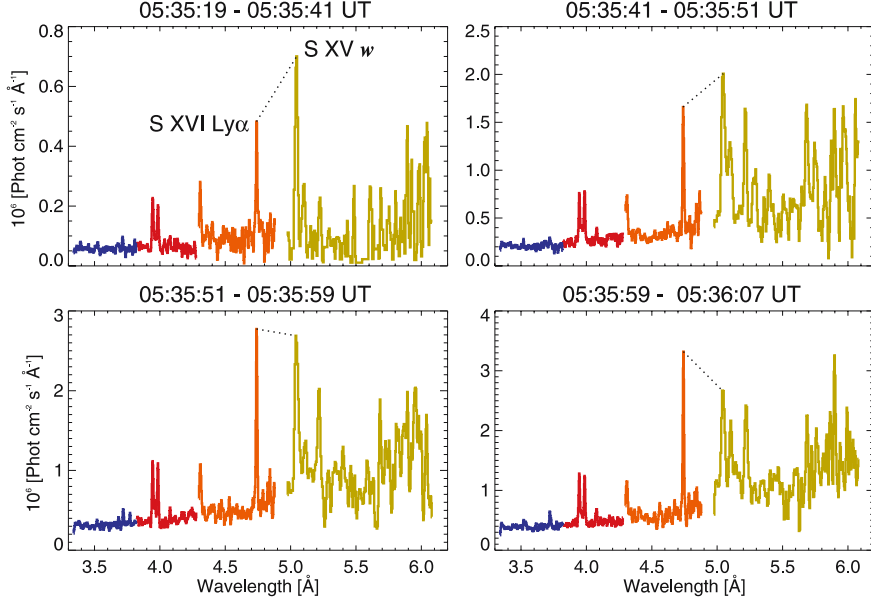


Figure 1. The four RESIK spectra observed during the rise phase of the flare on 2002 October 4. The data gathering intervals of consecutive spectra are 22, 10, 8 and 8 s respectively. Individual channels are presented in different colours.

Sylwester *et al.* (2003, 2005) and Kepa *et al.* (2006). In the figure two lines are labeled and connected with the dotted line in order to show a very fast rate of change in this wavelength band. Denoted lines correspond to the $1s-2p$ transitions in H- and He-like ions of S (at 4.73 Å and 5.02 Å respectively). The time variability of individual lines and/or bands observed by RESIK may be used as the basis for important plasma parameters determination. We have used the total fluxes (lines and continuum) measured in the most energetically separated RESIK channels (1 and 4) in order to study the average T_e and $EM = N_e^2 V$. These bands cover the 3.4 Å–3.8 Å and 5.0 Å–6.05 Å ranges respectively. In the left panel of Fig. 2, selected normalized light curves characterizing the different spectral regions are shown. As we have the volume V of the emitting region available (from reconstructed *RHESSI* images) we can derive the density N_e

$$N_e = \left(\frac{EM}{V} \right)^{\frac{1}{2}}, \quad (1)$$

and corresponding thermal energy content E_{th} of the region

$$E_{th} = 3k T_e (EM/V)^{\frac{1}{2}}. \quad (2)$$

Here k is the Boltzmann constant.

We have independently performed the corresponding calculations using *GOES* and RESIK data in two energy bands and the resulting thermal energy content is presented in the right panel of Fig. 2. It is seen that both instruments give very similar quantitative results with the maximum value of thermal energy content $E_{th} = 1.5 \times 10^{30}$ ergs.

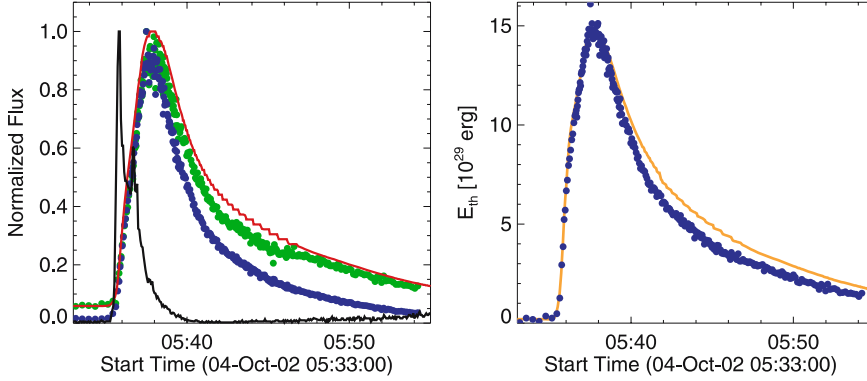


Figure 2. **Left panel:** The normalized light curves: thin black line corresponds to *RHESSI* data (25–50 keV), thin red to *GOES* 1–8 Å and two dotted lines (blue and green) correspond to RESIK channels’ 1 and 4 fluxes respectively. **Right panel:** The thermal energy content of the flaring region as obtained based on the *GOES* (thin red curve) and RESIK (thick blue dots) data.

Having had the fluxes in many spectral lines/bands observed instantaneously by RESIK we can calculate the momentary distributions of the plasma with the temperature (differential emission measure distributions, DEM) for particular times. In this respect, we have selected 18 narrow spectral bands containing intense lines and continuum from the RESIK spectra. The fluxes measured within these spectral bands have been time-integrated over a sequence of one minute intervals during the rise and the maximum phase of the event studied. Towards the decay longer intervals constituted the input set for the DEM calculations. The Withbroe-Sylwester (Sylwester *et al.* 1980) iterative DEM reconstruction algorithm has been used. Respective theoretical fluxes have been calculated using the CHIANTI code with the new coronal abundances (Sylwester *et al.* 2008) and Mazzotta ionization equilibrium adopted (Mazzotta *et al.* 1998). The uncertainties of DEM have been obtained from 100 Monte Carlo simulations of DEM calculations. The result of DEM calculations for one particular time interval centered at 05:49:05 UT is shown together with the uncertainties in the left panel of Fig. 3. The two-component nature of the distribution can be recognized easily. In the right panel the time evolution of DEM is shown in other representations. The two-component nature of DEM is preserved throughout the flare but the temperature and amount of plasma in each component is changing.

3. Concluding remarks

Although the assumption of isothermal approximation is oversimplified, using it with caution allows us to draw some important conclusions about the relative variability of temperature, thermal energy content and of plasma heating/cooling (see Jakimiec *et al.* 1992) evolution. For various times during flare evolution the DEM calculations indicate the presence of a two-component model: rather cool component (temperatures between 3 MK and 5 MK) and hot (temperatures between 9 MK and 18 MK). The relative amount of hot and cool plasma is changing during the evolution. The cool component is rather stable with regard to the amount of plasma and the peak temperature while the

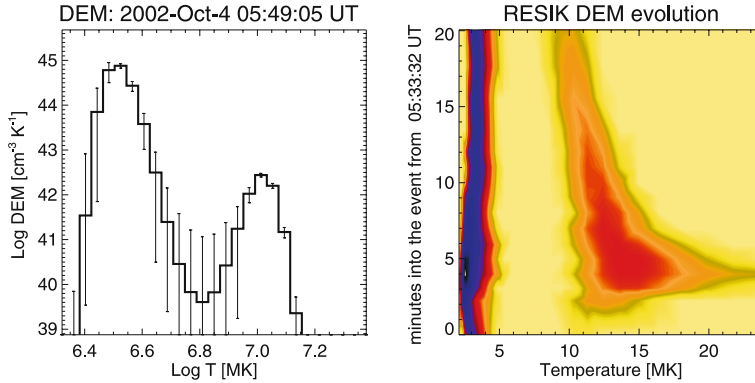


Figure 3. **Left panel:** The DEM distribution for one particular time with respective error bars indicated. **Right panel:** The stack (time sequence) of DEM shown in colour shade representation. The time covered (vertical axis) represents the flare evolution from the rise phase (bottom), through the maximum (around minutes 4–6) to the decay phase (top).

properties of hot component are very dynamic. During the cooling phase the maximum temperature of hot plasma component and its amount keep decreasing.

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Models of Solar Irradiance Variations: Current Status

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Abstract. Regular monitoring of solar irradiance has been carried out since 1978 to show that solar total and spectral irradiance varies at different time scales. Whereas variations on time scales of minutes to hours are due to solar oscillations and granulation, variations on longer time scales are driven by the evolution of the solar surface magnetic field. Here the most recent advances in modelling of solar irradiance variations on time scales longer than a day are briefly reviewed.

Key words. Sun: activity, faculae, plages, magnetic fields—solar irradiance—solar-terrestrial relations—sunspots.

1. Introduction

Solar irradiance is the total amount of solar energy at a given wavelength received at the top of the earth's atmosphere per unit time. When integrated over all wavelengths, this quantity is called the total solar irradiance, previously known as the solar constant. Since 1978 we know, however, that the Sun is a variable star and solar irradiance varies on all time scales at which it has been measured, i.e., minutes to decades (Willson & Hudson 1988, 1991; Fröhlich 2005), and most probably, on longer time scales as well (cf. Lockwood *et al.* 1999; Solanki *et al.* 2000).

This has important implications for our understanding of global changes in the earth's climate system. As weak as they are ($\approx 0.1\%$ over the solar cycle), irradiance variations affect the terrestrial environment (e.g., Eddy 1976; Reid 1987; Haigh 1996, 2001; Bond *et al.* 2001; Rind 2002; Langematz *et al.* 2005). Thus in order to assess quantitatively contributions of different sources of climate fluctuations, first the solar signal has to be reliably isolated. However, global climate changes typically occur on time scales longer than the length of the available irradiance record. This calls for a reconstruction of irradiance variations back to the pre-satellite period over as long a period as possible, which is only possible with the help of suitable models. Such models should include as much of our understanding of the physical mechanisms responsible for irradiance variations as possible and, to be successful, they must reproduce all available observations. Therefore, in the first part of this review (section 2) we concentrate on irradiance models aimed at recreating the measured variations in total (TSI) and spectral (SSI) solar irradiance. Then, in section 3, reconstructions of solar irradiance for earlier times are discussed.

2. Solar irradiance in cycles 21–23

2.1 Total solar irradiance

Solar irradiance variations on time scales of minutes to hours are caused by solar oscillations and granulation (Solanki *et al.* 2003). Significant changes in the Sun's luminosity and radius on time scales of 10^6 – 10^{10} years are predicted by the theory of stellar evolution. Variations on other time scales are predominantly due to the evolution of the solar surface magnetic field.

The two most striking features of the observed record of solar irradiance are the variations by about 0.1% in phase with the solar cycle and sharp dips with a comparable or even greater amplitude typically lasting 7 to 10 days. These dips are caused by the passage of sunspot groups across the visible solar disc as the Sun rotates. Towards activity maxima, when the number of sunspots grows dramatically, the frequency and depth of the dips increase. On average, however, the Sun brightens. This is due to the increased amounts of bright features, faculae and network elements, on the solar surface. The total area of the solar surface covered by such features rises much more strongly as the cycle progresses than the area of dark sunspots.

The most direct way to replicate the observed irradiance variations is thus to combine proxies (two or more) describing sunspot darkening (e.g., sunspot number or areas) and facular brightening (such as facular areas, Ca II or Mg II indices), e.g., using regressions (e.g., Chapman *et al.* 1996; Fröhlich & Lean 1998; Preminger *et al.* 2002; Jain & Hasan 2004). More physics-based models have also been developed (e.g., Fontenla *et al.* 1999, 2004; Fligge *et al.* 2000; Krivova *et al.* 2003; Wenzler *et al.* 2004, 2005, 2006; Giorgi *et al.* 2007).

In these models, maps of a given proxy, e.g., Ca II K brightness (PSPT models: Ermolli *et al.* 2003; Fontenla *et al.* 2004; Harder *et al.* 2005; Giorgi *et al.* 2007) or magnetogram signal (SATIRE models: see Krivova & Solanki 2005a; Solanki *et al.* 2005, and references therein) are analysed in order to extract information about the distribution of the magnetic features (sunspot umbrae, penumbrae, faculae, plages, the network, etc.) on the Sun's surface and its evolution with time. They are then converted into irradiance using semi-empirical model atmospheres such as those by Fontenla *et al.* (1999, 2006) and Unruh *et al.* (1999).

Figure 1 compares the total solar irradiance in cycles 21–23 calculated by Wenzler *et al.* (2006) using the SATIRE model and the Kitt Peak NSO magnetograms and continuum images with direct TSI measurements (Fröhlich 2006). The correlation coefficient between the data and the model is 0.91 for the whole period 1978–2003 and 0.94 for the period since 1992, when the new magnetograph came into operation at Kitt Peak and the quality of the data improved significantly. Moreover, if space-based magnetograms and continuum images from the MDI instrument on SoHO are used (which are only available since 1996), the correlation coefficient improves further reaching 0.96 (Krivova *et al.* 2003). Note that the reconstruction by Wenzler *et al.* (2006) extends the available TSI record by about 4 years going back to 1974, although with significant gaps.

2.2 Spectral solar irradiance

The advantage of the irradiance models involving model atmospheres is that both TSI and SSI can be calculated (Fontenla *et al.* 2004; Fontenla & Harder 2005;

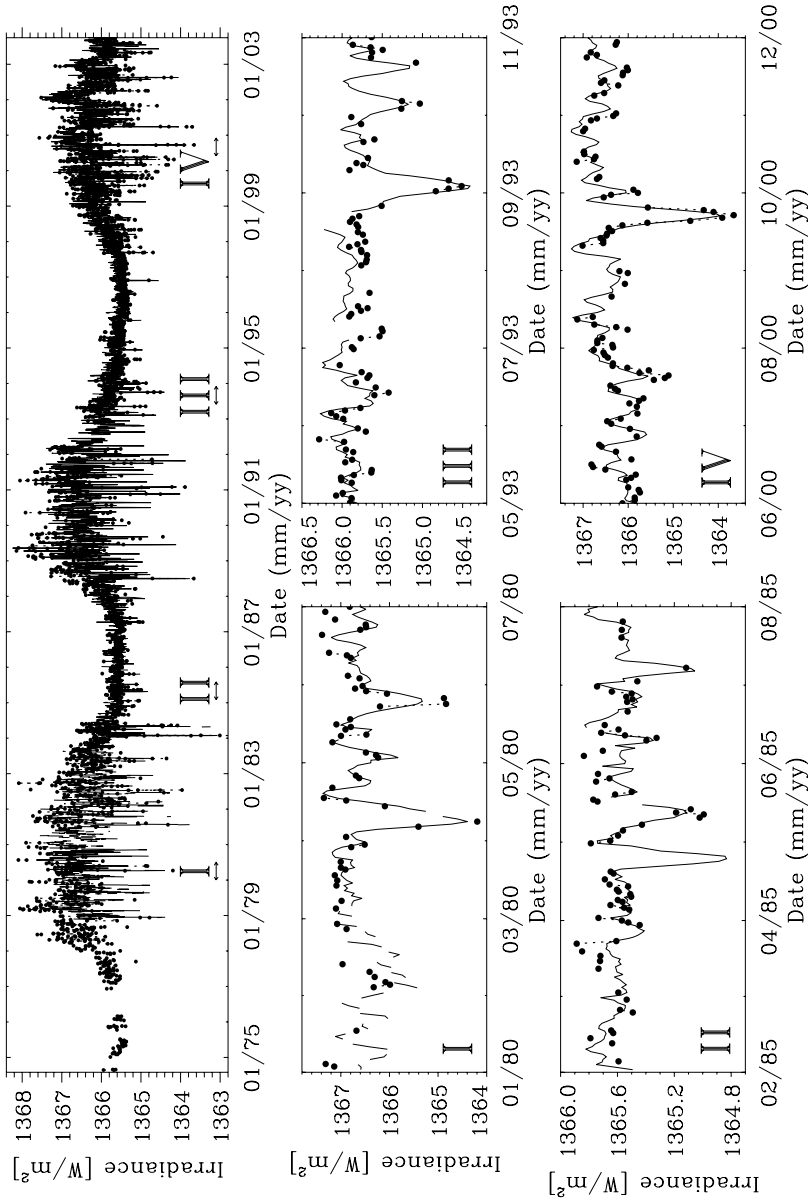


Figure 1. Top panel: The reconstructed TSI (filled circles, connected by dotted curve when there are no data gaps) based on NSO/KP data between 1974 and 2003. The solid line represents the measured total solar irradiance (PMOD composite, Fröhlich 2006) between 1978 and 2003. The **bottom panels** are enlargements of four shorter intervals at different activity levels from different cycles (from Wenzler *et al.* 2006).

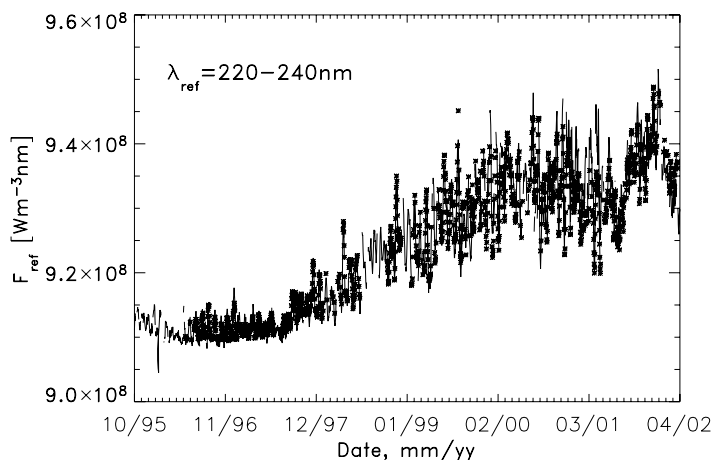


Figure 2. The solar irradiance integrated over the wavelength range 220–240 nm as a function of time. The solid line represents the SUSIM measurements (Floyd *et al.* 2003) and asterisks the SATIRE model. The correlation coefficient between the calculated and observed records is 0.97. (From Krivova *et al.* 2006).

Krivova & Solanki 2005b; Krivova *et al.* 2006; Danilovic *et al.* 2007). The models reproduce reasonably the observed SSI variations in the near-UV (above approximately 200 nm), visible and near-IR (Unruh *et al.* 1999; Krivova *et al.* 2003; Harder *et al.* 2005; Krivova & Solanki 2005b; Danilovic *et al.* 2007). One example is shown in Fig. 2, where the irradiance integrated over the wavelength range 220–240 nm obtained with SATIRE (asterisks) is compared to UARS SUSIM measurements. The correlation coefficient between the two time series is 0.97.

These models, however, fail at shorter wavelengths due to the LTE approximation which is often taken to compute model atmospheres but is no longer good in this spectral region. Employment of a non-LTE approximation is one possible line of attack on the problem (e.g., Fontenla *et al.* 1999, 2006). But since such calculations are much more arduous, non-LTE models remain as yet imperfectly developed and further efforts are needed in this direction (e.g., Haberreiter *et al.* 2005).

An alternative approach has recently been developed by Krivova & Solanki (2005b) and Krivova *et al.* (2006) that allows an empirical extrapolation of the SATIRE models down to 115 nm using available SUSIM measurements. Since the SATIRE model has been found to work well in the spectral range 220–240 nm (see Fig. 2), they used empirical relationships between the irradiance in this range and irradiance at other wavelengths covered by the SUSIM detectors (115–410 nm), to also calculate SSI at all these wavelengths. As an example, the calculated solar Ly- α irradiance is compared in Fig. 3 to SUSIM measurements and to the composite of UARS SOLSTICE measurements and related proxy models by Woods *et al.* (2000). Note that the latter record is adjusted to the UARS SOLSTICE absolute values, so that the difference between the two models is only due to the difference of about 5% in the magnitude of the Ly- α cycle shown by SUSIM and SOLSTICE. Other than that, the model agrees well with the completely independent

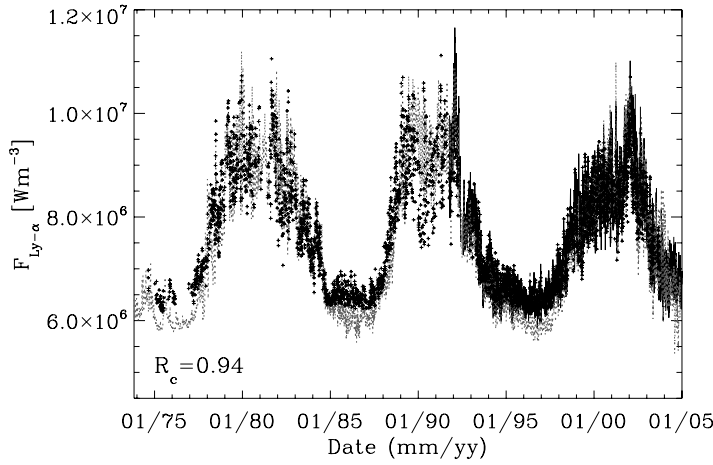


Figure 3. Solar Ly- α irradiance since 1974: reconstructed (plusses), measured by SUSIM (solid line) and compiled by Woods *et al.* (2000, dotted line). The correlation coefficient between the model and the composite is 0.94.

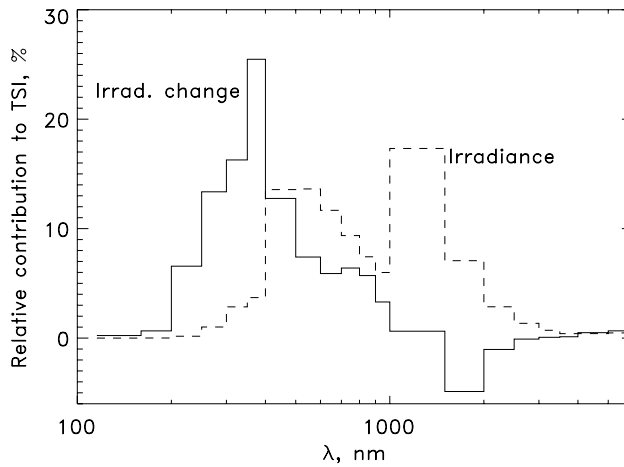


Figure 4. The relative contribution of different wavelength ranges to the TSI (dashed histogram) and its solar cycle variations (solid histogram). Note different sizes of bins: about 40 nm below 200 nm, 50 nm between 200 nm and 400 nm, 100 nm between 400 and 1000 nm and 500 nm at yet longer wavelengths. From Krivova *et al.* (2006).

composite time series, with a correlation coefficient of 0.94 (Krivova *et al.* in prep.).

Figure 4 shows the contribution of different spectral ranges to solar irradiance (dashed) and its solar cycle variation (solid) calculated by Krivova *et al.* (2006). The contribution of the UV irradiance variations to the total irradiance change is very high: about 60% of all TSI variations originate below 400 nm.

3. Going back in time

3.1 Solar irradiance since 1610

The remarkable agreement of the modelled variations of TSI and SSI with their measurements implies that it is the solar surface magnetic field which is responsible for their modulation. Thus the reconstruction of solar irradiance on longer time scales requires the computation of the Sun's magnetic field for the same time scale. When doing this, it is necessary to distinguish between cyclic variations, which can often be well described by proxies (e.g., the sunspot number), and secular variations, which are more tricky to estimate. A simple physical model has been developed by Solanki *et al.* (2000, 2002) that allows a reconstruction of the solar total and open magnetic flux from the sunspot number and provides an explanation for the origin of the secular variation. In this model, the sunspot number is used as a proxy for the freshly emerging magnetic flux in active regions. Emergence of magnetic flux in smaller ephemeral regions is considered separately. In doing so, it is taken into account that, firstly, the ephemeral regions start to emerge earlier than the active regions of the same activity cycle (e.g., Harvey 1992), and secondly, that the open flux lives much longer on the Sun's surface than the corresponding closed flux, e.g., in active regions. This leads to an overlap between activity cycles of the ephemeral regions and, thus to a build-up of the background magnetic flux, whose level varies with time.

This model has recently been used by Krivova *et al.* (2007) to reconstruct solar total irradiance back to 1700 using the Zurich sunspot number and to 1610 using the Group sunspot number. The reconstructed TSI is presented in Fig. 5. The model predicts an increase in the cycle-averaged TSI since the Maunder minimum of $1.3^{+0.2}_{-0.4} \text{ W m}^{-2}$, which is in agreement with other recent estimates. Wang *et al.* (2005) obtained a value of 1 W m^{-2} , but since they neglected the extended length of the ephemeral region cycle, this should be considered as a lower limit. On the other hand, Foster (2004), cf. Lockwood (2005), obtained an upper limit of 1.7 W m^{-2} by assuming that the network disappeared during the Maunder minimum.

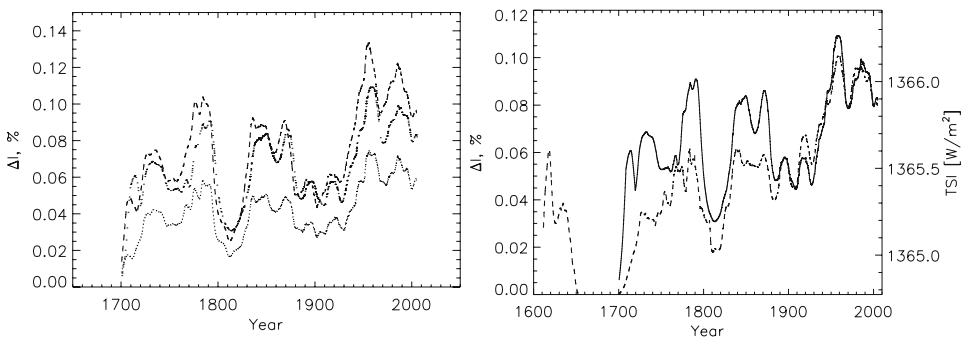


Figure 5. **Left:** 11-yr running mean of the reconstructed TSI based on the Zurich sunspot number since 1700 (solid line). The dashed and dotted lines show the models obtained under extreme assumptions and represent the upper and lower limits on the magnitude of the secular increase in the TSI, respectively. **Right:** Same but now for the TSI reconstructed using Zurich (solid line) and Group number (dashed line). (From Krivova *et al.* 2007).

3.2 Solar irradiance before 1610

On even longer time scales no direct measurements of solar variability are available. Therefore, indirect proxies must be used, such as the cosmogenic isotopes ^{10}Be and ^{14}C , which are produced when high-energy cosmic rays enter the earth's atmosphere and react with nitrogen and other atoms. Since the cosmic ray flux at the earth is modulated by the Sun's magnetic activity, it is possible to reconstruct the strength of the solar open magnetic flux using the production rates of cosmogenic isotopes stored in natural terrestrial archives. This allows a reconstruction of the sunspot number (Solanki *et al.* 2004; Usoskin *et al.* 2006) over the last 7 millennia (or even 11, although somewhat less reliably). This reconstruction has shown that the Sun is currently in a state of unusually high magnetic activity.

Since only a cycle-averaged sunspot number can be obtained, the model described in section 3.1 cannot be applied in a straightforward manner to calculate the variations of the TSI. Nevertheless, analysing the behaviour of the cycle-averaged sunspot number, solar magnetic flux and the TSI during the last 4 centuries and applying these relationships to the complete reconstructed time series of the sunspot number, it is possible to assess the evolution of the solar irradiance over the last few millennia. Such a reconstruction is currently on the way (Vieira *et al.* in prep.).

4. Summary and outlook

Models of solar irradiance have progressed significantly in the past years. Observed variations of solar total and spectral irradiance are reproduced with high accuracy. Advances have also been made in reconstructing the past behaviour of the solar irradiance, on time scales of centuries to millennia. It has been shown that solar total irradiance has increased by approximately 0.1% since the end of the Maunder minimum, which is comparable to the 11-yr modulation during the 3 recent activity cycles but is significantly less than was previously assumed. On even longer time scales, the Sun appears to be in an unusually active state: over a period of 11,400 years, a similar level of activity as in the last 60 years was only reached during about 3% of the time.

Despite this progress, there is still considerable work to be done. The following developments are needed, some of which can be expected in the near future:

- removal of the remaining free parameter(s) from the models;
- reconstruction of the solar spectral irradiance back to the Maunder minimum;
- reconstruction of the TSI and SSI over the last 7000 years;
- improvement of the employed semiempirical model atmospheres using new sunspot and facular observations (e.g., Ortiz *et al.* 2002; Ermolli *et al.* 2007; Mathew *et al.* 2006) and measurements of SSI from SORCE and SCHIAMACHY;
- non-LTE model atmospheres.

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Spatially Resolved Images and Solar Irradiance Variability

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Abstract. The Sun is the primary source of energy that governs both the terrestrial climate and near-earth space environment. Variations in UV irradiances seen at earth are the sum of global (solar dynamo) to regional (active region, plage, network, bright points and background) solar magnetic activities that can be identified through spatially resolved photospheric, chromospheric and coronal features. In this research, the images of CaII K-line (NSO/Sac Peak) have been analysed to segregate the various chromospheric features. We derived the different indices and estimated their contribution from the time series data to total CaII K emission flux and UV irradiance variability. A part of the important results from this research is discussed in this paper.

Key words. Solar chromosphere: spatially resolved images—chromospheric features—irradiance variability.

1. Introduction

Since the radiative output of the Sun is one of the main driving forces of the terrestrial atmosphere and climate system, the study of the changes in the solar energy flux has become of interest. Although the long-term change in total solar irradiance (the solar energy flux integrated over the entire spectrum) can be considered to be one of the major natural forces of the earth's climate system, the study of UV irradiance variability is an equally important issue in solar physics. The UV irradiance variability has profound influence on the earth's climate and space weather. For more than 2 to 3 decades, the solar irradiance (both bolometric and at various wavelengths) has been monitored as 'Sun as a star' from several satellites. It is known that the solar energy flux changes over a solar cycle. It has been shown that the long-term irradiance variations are attributed to the changing emission of bright magnetic elements (Foukal & Lean 1988; Kariyappa & Pap 1996, 2000; Worden *et al.* 1998), whereas the short-term irradiance variations are directly associated with active regions as they evolve and move across the solar disk (Lean 1987). The difference between the observed and model irradiance variability has not been explained and further investigations are required. The current irradiance models are based on full-disk integrated flux, such as radio flux, CaII K 1 A index, HeI equivalent width, but may not be from spatially resolved features of the Sun. In order to identify and understand the underlying physical mechanisms of solar irradiance variability and to estimate the contribution of various chromospheric features to UV irradiance, a detailed analysis of spatially resolved data both from ground

and space are required. The CaII H and K resonance lines are widely used to study the solar chromospheric structures and these lines are very sensitive to the variations in temperature and the magnetic field strength. Therefore they are excellent indicators of the chromospheric structural changes related to solar magnetic activity. The analysis of the CaII H and K spectro-heliograms demonstrate that the main features which are responsible for chromospheric emission are the plages, the network, internetwork and background chromospheric regions. However, separating out these features from the images is one of the biggest problems in image analysis.

2. Observations and data analysis

The CaII K spectro-heliograms have been obtained from NSO/Sac Peak using the spectro-heliograph. We have selected and analyzed the daily images of 6 years (namely 1980, 1985, 1987, 1988, 1989 and 1992) representing the different phases of the solar activity. We derived the full disk indices (spatial K index and Full Width at Half Maximum, FWHM, of the histogram taken for the full disk image), the quiet Sun intensity, the cumulative intensity and area of the individual chromospheric features using the histogram plot taken for the full-disk image of the chromosphere. In this paper, we discuss the variations of these indices and estimate the contribution of the various chromospheric features to CaII K & UV irradiance variability.

3. Results and discussion

We have calculated the histograms for the full-disk images (Kariyappa & Pap 1996, 1999; Kariyappa 2000) to analyze the CaII K spectro-heliograms and to separate out the various chromospheric-magnetic features. We have used the morphology and intensity level of the features as the main criteria to segregate the different features from the images. We are familiar with the morphology of different features:

- the plages are very bright, large, and compact structures,
- the network elements are cellular structures with bright boundaries, and
- the remaining features relate to background + internetwork regions.

We first assumed the intensity levels that might bound the plage pixels in a histogram, and then examined images in which the pixels with greater or lesser intensities than those bounds were masked. The bounding intensities were then adjusted until the masked images mapped the plage regions. We applied similar techniques for the network and internetwork + background features. The intensity and the number of pixels (area) for plages, network, and internetwork + background regions have been derived from these images using their corresponding histogram plots. It has been established that the spatial K index, the FWHM, the intensity of different chromospheric features are well correlated with UV irradiance variability measured in MgII h and k lines and whereas the area of the chromospheric network elements is anticorrelated with solar activity cycle (Kariyappa & Sivaraman 1994; Kariyappa & Pap 1996; Kariyappa 2000). The anticorrelation of the area of the network with solar cycle has also been shown by others (Muller & Roudier 1994; Ermolli *et al.* 1998; Berrilli *et al.* 1998, 1999). The network is fainter during solar minimum, but covers a larger area on the solar disk and hence their contribution should be considered in constructing the irradiance models.

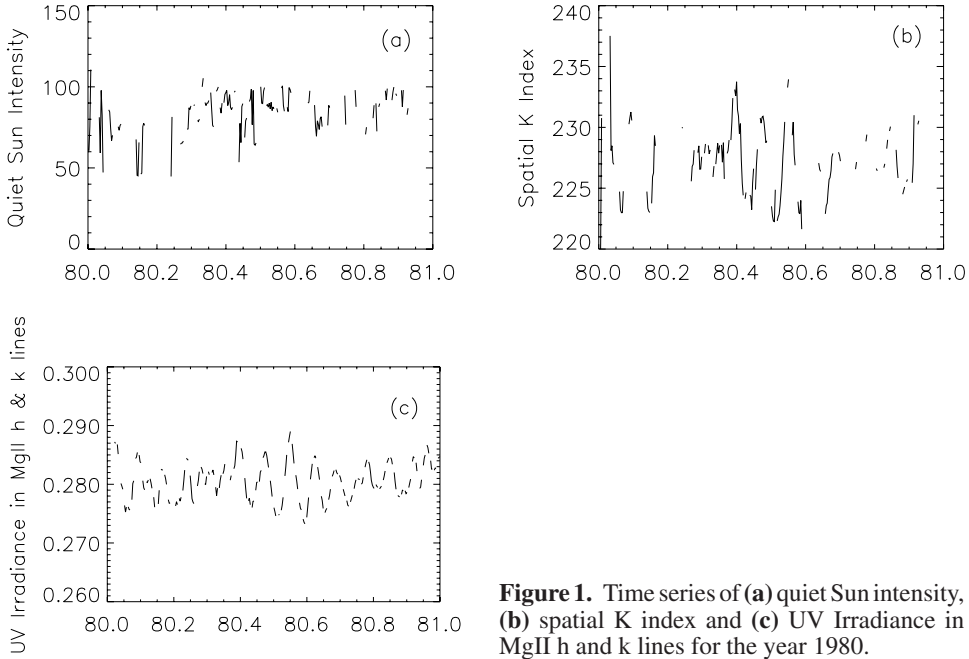


Figure 1. Time series of (a) quiet Sun intensity, (b) spatial K index and (c) UV Irradiance in MgII h and k lines for the year 1980.

In addition to full-disk indices and intensity and area of plages, network and internetwork + background regions, we derived the quiet Sun intensity in a very quiet region at the center of the solar disk and plotted them as a function of time. In Fig. 1, we have compared the variation of quiet Sun intensity with spatial K index and UV irradiance measured in MgII h and k lines for the year 1980 and found that they are well correlated with each other. This represents that quiet Sun intensity will vary in phase with the solar activity and will contribute to the variation of UV irradiance and its contribution should be considered in irradiance models. The results show that the intensity of the quiet Sun will be varying with the solar cycle and is very dynamic and is crucial for the understanding of the chromosphere itself.

From the intensity time series of individual features, we estimated the contribution of various chromospheric features to CaII K irradiance variability using the following expression:

$$\text{Contribution (in \%)} = (\text{Intensity of chromospheric feature} / \text{spatial K index}) \times 100,$$

where the spatial K index is the full-disk intensity.

We derived the average contribution for the whole year and found that about 50% of the CaII K solar cycle variability results from plage, about 32% from the network, and about 18% combined from internetwork + background regions. The percentage of contributions of the three features will also explain the UV irradiance variability measured in MgII h and k lines, since it has been established that there is a good correlation between the variability of CaII K emission and UV irradiance. We can conclude from these results that besides the plages, a significant portion of the variations observed in UV irradiance are related to the changing emission of the network and

internetwork + background regions. They will play a significant role in their contribution to the variation in UV irradiance and are not constant as assumed in the irradiance models.

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Wave Heating of the Solar Chromosphere

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Abstract. The nonmagnetic interior of supergranulation cells has been thought since the 1940s to be heated by the dissipation of acoustic waves. But all attempts to measure the acoustic flux have failed to show sufficient energy for chromospheric heating. Recent space observations with TRACE, for example, have found 10% or less of the necessary flux. To explain the missing energy it has been speculated that the nonmagnetic chromosphere is heated mainly by waves related to the magnetic field. If that were correct, the whole chromosphere, magnetic as well as nonmagnetic, would be heated mainly by waves related to the magnetic field. But contrary to expectation, the radiation emerging from the nonmagnetic chromosphere shows none of the signatures of magnetic waves, only those of acoustic waves. Nearly all the heating of the nonmagnetic chromosphere must therefore be due to acoustic waves. In the magnetic network on the boundary of supergranulation cells, on the other hand, the small filling factor of the magnetic field in the photosphere implies that only a small fraction of the wave flux that travels upward to heat the chromosphere can be channeled by the magnetic field. Hence, while some of the energy that is dissipated in the magnetic network is in the form of magnetic waves, most of it must be in the form of acoustic waves. Thus, the quiet solar chromosphere, instead of being heated mainly by magnetic waves throughout, must be heated mainly by acoustic waves throughout. The full wave flux heating the quiet chromosphere must travel through the photosphere. In the nonmagnetic medium, this flux is essentially all in the form of acoustic waves; TRACE registers at most 10% of it, perhaps because of limited spatial resolution.

Key words. Sun: chromosphere, oscillations—heating and dynamics—acoustic waves—magnetic waves.

1. Introduction

The observation that the solar chromosphere is hotter than the photosphere implies a heating mechanism that is different from radiative and convective. Biermann (1946) and Schwarzschild (1948) recognized that the heat source had to be the turbulent convection which generates acoustic waves whose dissipation provides the energy for chromospheric heating. The paper is structured as follows: In section 2, we consider

measurements of the acoustic flux heating the chromosphere and describe radiative signatures of acoustic and magnetic waves and in section 3, we investigate the distribution of magnetic fields in the quiet chromosphere.

2. Observations of the acoustic flux

An understanding of the structure of the nonmagnetic chromosphere, as described by the models of Vernazza *et al.* (1981; hereafter VAL), requires a measurement of the energy flux of acoustic waves, as well as their velocity spectrum, and construction of a model. Observations of the acoustic spectrum are difficult because of limitations of both ground-based as well as space-based measurements as regards spatial and temporal resolution.

The energy flux in acoustic waves observed by Deubner and reported by Ulmschneider (1990) showed that at a height of 900 km, where the radiative cooling function per unit volume of VAL model C has its maximum, the wave flux is deficient by an order of magnitude. Wunnenberg *et al.* (2002) found that the wave sources which their observations imply lie preferentially above intergranular lanes; they are intermittent also in time. Both features are inconsistent with our understanding of continuous wave emission from the convection zone. Fossum & Carlsson (2005) observed intensity fluctuations from space with TRACE in the 1600 Å passband and found that the waves contained at most 10% of the energy flux required for chromospheric heating. They concluded that the nonmagnetic chromosphere must therefore be heated by waves related to the magnetic field (cf. Carlsson 2007). This idea can be tested by investigating signatures of the nature of waves seen in the emergent radiation from the chromosphere.

Emission from the chromosphere reveals the nature of the waves producing the radiation, in the asymmetry and evolution of the intensity, the dominant periods of fluctuations, and the magnitude of the average intensity.

In the magnetic network, the intensity in the H and K lines is more or less symmetric (Fig. 1 of Lites *et al.* 1993), with mainly simultaneous blue and red peak enhancements (frequency of occurrence in the K line of 80%, Grossmann-Doerth *et al.* 1974) and in the internetwork chromosphere, with mainly single blue-peak enhancements (frequency of 40%).

A second indicator of the nature of the waves is the time dependence of the line intensity. Figure 2 of Lites *et al.* (1993) shows the H line as a function of time and wavelength near line center; the inner panels are for the intensity of a network bright point, the outer panels, for an internetwork H_{2v} bright point. The network intensity is high most of the time and exhibits a high degree of symmetry; the intensity from oscillations in the internetwork medium is sharply peaked in time on the blue side of the line center, with the peak intensity dropping to the level of the background in a time interval of 5% of the 3 min wave period. In addition, there are long gaps in the oscillations when the emergent radiation shows only the background.

A third indicator of the nature of a wave source is the power spectrum of velocity variation, seen for the H line in Fig. 6 of Lites *et al.* (1993) and for an infrared line of Ca II in Fig. 2 of Deubner & Fleck (1990); whereas power in the network is high at low frequencies of 2–3 mHz, and low at higher frequencies of 5 mHz, power in the internetwork medium is low at low frequencies and high at higher frequencies.

A fourth indicator is seen in the filtergram (Fig. 5, from von Uexküll & Kneer 1995) exhibiting fluctuations in the K_{2v} emission peak. At the locations in the internetwork chromosphere, which was identified by a low time-average intensity, the maximal intensity exceeds the cell-interior average by a factor of at least 1.5. Von Uexküll and Kneer have identified time intervals ranging from 150 s to 188 s for fluctuations covering at least two adjacent intervals. These oscillations are clearly of the 3 min type of acoustic oscillations. No long-period oscillations are in evidence. The straightforward interpretation of these fluctuations is that they are associated with acoustic waves, and that waves related to magnetic fields are not involved.

3. Topology of the quiet chromosphere

The emission in the H and K lines separates stars with outer convection zones broadly into two groups (Noyes *et al.* 1984), one with high emission from rapidly-rotating young stars, and the other with lower emission from slowly-rotating old stars. The two groups differ by two to three orders of magnitude in the power of the chromospheric emission in the cores of the H and K lines. The high emission of fast rotators is thought to represent the effect of magnetic activity whereas the low emission of slow rotators represents the effect of basal acoustic heating (Schrijver *et al.* 1989).

The Sun is a slowly rotating main sequence star. The bulk of the nonradiative heating is due to acoustic waves. But the intensity enhancement in the K_{2v} emission of bright points in the magnetic network over that in the dark supergranulation cell interior is only 27% (Skumanich *et al.* 1975). With a filling factor of the network in the K line of 39%, the net increase of the emission from the quiet Sun over that of the dark cell interior alone is only 10%, which implies that even in the magnetic network a large part of the emission is due to acoustic waves.

The filling factor of the network of 39% at a height of 1000 km allows us to infer the gross topology of the magnetic field, assuming validity of the equations for a slender flux tube. The filling factor $f(z)$, of the magnetic field as a function of height z , then varies on a length scale of twice the density scale height, $\mathcal{H}$, i.e., $f(z) \propto \exp((z - z_0)/2\mathcal{H})$. With $\mathcal{H} \sim 100$ km, the filling factor at the top of the photosphere is $f(500 \text{ km}) \sim 3\%$, and in the middle, $f(250 \text{ km}) \sim 0.9\%$. For the idealized “thin” flux tube, the top of the nonmagnetic chromosphere, where $f \sim 100\%$, would be reached at 1200 km. The filling factor at $z \sim 250$ km in the photosphere allows us to infer broadly the properties of heating by magnetic waves in the network. The velocity amplitude of acoustic waves in Ca II bright points in the observations of Lites *et al.* (1993) has a Mach number of about 10% (Carlsson, private communication). Since the velocity amplitude of flux tube waves varies with a length scale of $4 \times \mathcal{H}$, the Mach number increases to about 20% at the base of the chromosphere ($z = 500$ km). If the initial velocity amplitude was much higher than it is at internetwork bright points, the limiting value of the Mach number of 0.3 for dissipation (Schwarzschild 1948; Hasan *et al.* 2003) would be reached in the photosphere and dissipation would set in, without significantly increasing the energy flux to the chromosphere. Thus, magnetic waves in the magnetic network can contribute at most a few per cent to chromospheric heating. We conclude that all the heating in the internetwork chromosphere is due to acoustic waves, as is most of the heating in the magnetic network.

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Ultraviolet Spectroscopic Observations of Coronal Streamers in the SOHO Era

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Abstract. Measurements made with the Ultraviolet Coronagraph Spectrometer (UVCS) on the Solar and Heliospheric Observatory can be used to determine physical parameters in the solar corona such as hydrogen and ion kinetic temperatures, electron densities, and absolute elemental abundances. Hydrogen and ion outflow velocities can be determined by combining the UV spectroscopic measurements with white light polarized brightness measurements. These combined measurements can be used to reveal physical characteristics of coronal streamers. To date we have studied plasma properties, such as the variation of plasma outflows in quiescent streamers, primarily in classic helmet streamers at solar minimum. Outflows have not been observed in the centers of coronal streamers suggesting that these are closed magnetic field regions. We propose to study all of the coronal streamers in the UVCS synoptic dataset in order to investigate different types of streamers and their long-term evolution.

Key words. Sun—corona—ultraviolet.

1. Introduction

Coronal streamers are bright ray-like structures extending outward from the Sun for many tens of solar radii. These structures are bright because they are more dense than the background corona. Their higher densities are due to particle trapping by the large-scale coronal magnetic field. With the Ultraviolet Coronagraph Spectrometer (UVCS) onboard the Solar and Heliospheric Observatory (SOHO), we have made considerable progress in understanding the energetics, composition, and dynamics of streamers. In this paper we describe the key results from UVCS observations of coronal streamers and the progress being made on establishing a 10-year database of streamer properties covering most of Solar Cycle 23. The database properties are derived primarily from the UVCS synoptic observations.

The goals of the UVCS streamer study are to better understand the properties (densities, temperatures, abundances and outflow speeds) of coronal streamers and to relate these properties to characteristics of the solar wind. Important questions to be answered are the following:

- From which part of the coronal streamer does the solar wind originate?
- Do the same physical processes that occur in coronal holes operate in coronal streamers?

- What is the three dimensional structure of a typical coronal streamer?
- How do coronal streamer properties vary with the solar cycle?
- How do changes in coronal streamers relate to changes in the solar wind?

2. Results from UVCS streamer studies

The UVCS program (Kohl *et al.* 1997) has made important contributions to the understanding of coronal streamers. Some of the key findings are listed below:

1. Outflow was measured at the tips of streamers at high heights and also near the edges of streamers at low heights (Strachan *et al.* 2002). The outflow speeds are consistent with LASCO ‘blob’ speeds at higher heights.
2. Abundances for O, Mg, Al, Si, S, and Fe in streamer cores were found to be severely depleted compared to streamer legs (e.g., flanks), irrespective of the FIP effect. In addition, the similarities between the abundances (for these same ions) in the streamer legs and the solar wind suggest that the legs are the origin of the slow solar wind (Raymond *et al.* 1997).
3. UVCS observed that the O VI 103.2 nm emission in the streamer core was considerably less intense than in the legs. This is now believed to be due to a depletion in the oxygen elemental abundance relative to photospheric values (Noci *et al.* 1997; Raymond *et al.* 1997). This may be due to gravitational settling in closed field regions.
4. Additional studies showed that not all quiescent streamers have large minor ion abundance depletions in their cores. Active region streamers have a wider range of ion abundances in their cores compared to quiescent streamers (Uzzo *et al.* 2006).
5. O^{5+} kinetic temperatures for equatorial streamers at Solar Minimum are higher than the temperatures for Solar Maximum streamers at mid-to-high latitudes. Presently there is no definitive explanation for this trend although it could be a density effect (see Frazin *et al.* 1999; Strachan *et al.* 2004).
6. Extreme ion (O^{5+}) kinetic temperatures and temperature anisotropies ($T_{\parallel}$ vs. $T_{\perp}$) exist at high heights in quiescent streamers. This similarity to coronal hole anisotropies suggest that the same mechanism for plasma heating and acceleration (ion cyclotron resonance absorption) occurs in coronal holes as well as in streamers, although at higher heights in streamers (Frazin *et al.* 2003).

3. New streamer studies using UVCS synoptic data

UVCS has made ‘full corona’ maps in H I Ly α and the O VI doublet (103.2 and 103.7 nm) either daily or on alternating days for almost a complete solar cycle. These data can be used to track changes in the physical parameters of coronal streamers as a function of magnetic field strength and field line orientation as derived from extrapolated photospheric magnetic fields. The synoptic maps are made from 27 individual ‘full corona’ plane-of-the-sky maps built from spatial scans to cover heliocentric heights from 1.5 to 3 $R_{\odot}$ and 360° around the corona. Intensities and line widths at a fixed heliocentric height from each daily map are replotted on a longitude (time) vs. latitude grid to build Carrington rotation maps like those shown in Fig. 1. The maps in the figure are for 2 $R_{\odot}$ but other heights can be plotted as well to show streamer changes as a function of height. Carrington maps are useful for tracking the location of streamers

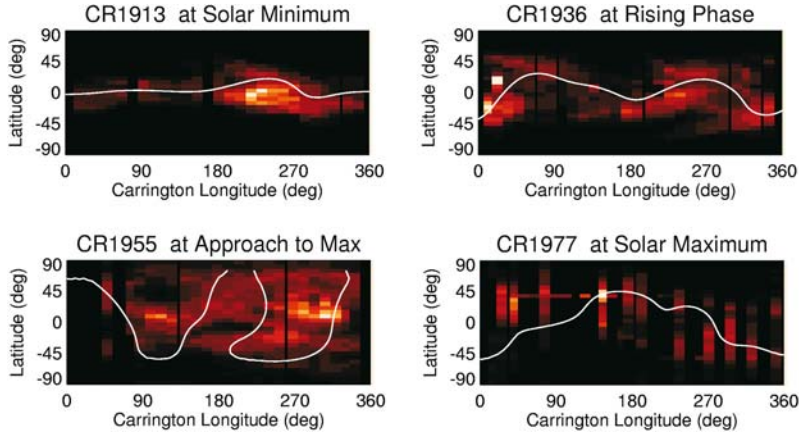


Figure 1. Representative O VI 103.2 intensity maps from different periods of solar cycle 23. White curves show the magnetic neutral line at $2.5 R_{\odot}$ from WSO data.

over many solar rotations. We use the magnetic neutral line plots from the Wilcox Solar Observatory (WSO) to guide the automatic selection of the most prominent streamers. These are the ‘edge-on’ streamers that appear when the magnetic neutral line is horizontal (i.e., the current sheet is perpendicular to the plane of the sky).

Representative data from different periods of the solar cycle are shown in Fig. 1. At solar minimum, streamers are formed mainly along the magnetic neutral line (heliospheric current sheet) but as the solar cycle progresses, the neutral line becomes warped and streamers form at all latitudes. Active region streamers have the largest deviations from the neutral line. Currently these are removed from the dataset but they will be examined as a separate category once we are finished with the quiescent streamers that have well-defined helmets (along the neutral line).

We are developing software which automatically detects ‘edge-on’ streamers in the UVCS synoptic dataset. The goal is to retrieve a subset of streamers, all with similar geometric configurations, so that *intrinsic* changes in the plasma parameters for streamers can be identified. The routines are used to extract H I Ly α and the O VI doublet line intensities, line ratios, and line widths from the 10+ years of UVCS synoptic data. From these data, quantities such as kinetic temperatures and outflow velocities can be derived. We will use these data to identify long term trends in streamer properties which cannot be obtained from observations of individual streamers.

4. Summary

UVCS spectroscopic observations have revolutionized our knowledge of the plasma state in coronal streamers, and have clarified the connections between streamers and the slow speed wind. While most studies have focused on individual streamers, the UVCS synoptic database will allow the entire range of streamer variability over the solar cycle to be studied. It is desirable to continue such long range studies into the new solar cycle with coronagraph instruments capable of imaging and spectroscopy over a larger instantaneous field of view.

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Microflares as Possible Sources for Coronal Heating

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Abstract. We present a preliminary study of 27 microflares observed by Solar X-ray Spectrometer (SOXS) mission during July 2003 to August 2006. We found that all 27 microflares show the Fe-line feature peaking around 6.7 keV, which is an indicator of the presence of coronal plasma temperature ≥ 9 MK. On the other hand, the spectra of microflares show hybrid model of thermal and non-thermal emission, which further supports them as possible sources of coronal heating. Our results based on the analysis show that the energy relapsed by the microflares is good enough for heating of the active corona. We discuss our results in the light of the hybrid model of microflares production.

Key words. Solar flares: microflares, Fe-line feature—coronal heating.

1. Introduction

The solar corona is some 200 times hotter than the surface of the Sun. The main source of energy for heating the corona is believed to be the magnetic field which dominates the corona. Magnetic reconnection is probably the most important mechanism for releasing magnetic energy and may, therefore, be important for coronal heating or micro-flaring. The best observational examples of reconnection in the corona are thought to be X-ray bright points, which are small-scale brightenings seen randomly throughout the whole corona. These brightenings are called as microflares and further small scale brightening as nanoflares. The coronal heating mechanisms so far widely known are based on two fundamental physical processes viz., acoustic wave dissipation and micro/nanoflares. Acoustic heating is ruled out due to the dependency of Lx on rotational speed. Magneto-acoustic heating by waves and/or particles induced from magnetic fields at the bottom of the corona is, however, proposed as a fundamental source. However, the observational fact that the Sun emits a large number of strong bursts of X-rays before the UV emission of flares, known as ‘nanoflares’ was proposed as a mechanism to heat the corona. Parker (1988) proposed that at 250,000 K some magnetic reconnection takes place between chromosphere and corona. Thus low altitude reconnection in the upper chromosphere or low transition region might be the possible source of energy release to heating the corona. The SOHO observed low temperature gas ($< 400,000$ K) going down; hotter gas going up and nowadays it is thought to be one of the mechanisms involved in coronal heating. In this context, one of the most potential candidates for coronal heating proposed in the last decade is micro and

nanoflares, which have been observed in EUV and X-ray waveband by EIT/SOHO and SXT/Yohkoh missions. It has been proposed that these small-scale flares exhibit hard X-ray spectrum indicative of electron acceleration process in them and these electrons move down and heat the lower corona and transition region layers. Recently it has been proposed that microflares show thermal plus non-thermal hybrid mechanism, which may indicate them as a strong source for coronal heating (Jain *et al.* 2006b).

The Solar X-ray Spectrometer (SOXS) mission (Jain *et al.* 2005, 2006a) observes full disk integrated X-ray emission from the Sun to observe solar flares of low intensity to very high intensity and has been found to be a very useful instrument to study microflares. We have considered the microflares from SOXS mission simultaneously to other optical and radio wavebands in order to confirm their occurrence and identification for the current investigation. The SOXS mission data provide excellent spectra in the energy range 4–56 keV using Si and CZT detectors. The X-ray spectra from Si detector in the energy range 4–25 keV, enable us to study Fe and Fe/Ni line features. The Fe and Fe/Ni X-ray emission features are not visible in the coronal X-ray emission due to gravitational settling unless the temperature exceeds 9 MK during flares (Jain *et al.* 2005, 2006a). 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 the small impulsive microflare events originate 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.

In the present paper, we have studied 27 microflares using temporal and spectral data from the Si detector of the SOXS mission. We aim to address the question whether microflares are the most possible candidates for coronal heating.

2. Observations

Flares on GOES intensity class $\leq C1.0$ are defined as microflares (Chang *et al.* 2004; Jain *et al.* 2006b). 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. Thus we have chosen only those SOXS flares as microflares, which have been observed by GOES on the intensity class $\leq C1.0$ and also observed either in $H\alpha$ or radio wavebands. The list of microflares observed by SOXS mission and chosen for the current investigation is presented in Table 1.

3. Analysis and results

Details of SOXS mission, in-flight performance, calibration, instrumental response and background are described by Jain *et al.* (2005, 2006a). The data format and data analysis techniques are also presented by them. SOXS delivers two types of data, viz., temporal and spectral mode. After unambiguous identification of the microflare, it is cross-correlated with GOES and $H\alpha$ or radio observations on time and intensity scales. Then we undertook detailed spectral analysis for the period of microflare using standard OSPEX software in SolarSoft package. Fig. 1 shows, as an example, the light

Table 1. Temporal characteristics of microflares.

Sl.no.	Date	GOES class	Start time			Peak time			End time			Duration		Peak counts/s
			Hour	Min	Sec	Hour	Min	Sec	Hour	Min	Sec	Min	Sec	
1	23 Dec 03	B8.7	4	55	7.1	4	57	3.2	5	1	48.9	6.70		48
2	11 Jan 04	C1.0	4	14	53.3	4	16	29.3	4	21	30.3	6.62		77
3	01 Feb 04	B6.9	6	43	33.8	6	49	36.8	6	52	7.9	8.57		68
4	25 Feb 04	B8.9	4	6	19.7	4	10	0.4	4	14	40.7	8.35		65
5	25 Feb 04	B9.8	4	59	20.5	5	6	56.2	5	24	36.8	25.27		104
6	25 Feb 04	C1.0	5	24	36.8	5	27	21.2	5	36	3.5	11.45		68
7	04 Mar 04	B4.6	4	49	4.3	4	53	56.7	5	12	41.6	23.62		113
8	06 Apr 04	C1.0	5	34	11.2	5	35	6.8	5	39	28.1	5.28		141
9	24 Oct 04	B7.1	3	46	18.4	3	48	10.5	3	55	45.6	9.45		31
10	24 Oct 04	B8.6	4	10	45.3	4	12	42.6	4	16	37.3	5.87		94
11	24 Oct 04	B9.5	4	26	31.9	4	28	17.6	4	31	5.8	4.57		93
12	24 Oct 04	B7.1	5	52	5.3	5	53	26.6	5	56	9.2	4.07		38
13	30 Oct 04	B9.2	4	52	0	4	56	0	5	1	0	9.00		176
14	30 Nov 04	B7.7	3	56	23.1	3	59	22.2	4	3	50.9	7.46		51

Table 1. (Continued)

Sl.no.	Date	GOES class	Start time			Peak time			End time			Duration		Peak counts/s
			Hour	Min	Sec	Hour	Min	Sec	Hour	Min	Sec	Min	Sec	
15	31 Jan 05	B5.9	6	17	20.7	6	18	24.3	6	22	2.4	4.70		70
16	06 Apr 05	B7.3	4	44	49.8	4	48	58.6	4	56	59.9	12.17		78
17	24 Jan 06	B3.8	4	10	20.9	4	12	36.3	4	16	52.3	6.52		37
18	24 Jan 06	B3.3	5	21	18.6	5	24	6.3	5	27	23.9	6.09		38
19	20 Mar 06	B7.1	4	35	30.2	4	37	3.2	4	40	15.1	4.75		63
20	11 Apr 06	C1.0	4	36	24.6	4	38	14.4	4	42	58.1	6.56		133
21	24 Apr 06	C1.0	3	58	37.7	4	2	0.6	4	14	25.9	15.80		104
22	26 Apr 06	B3.7	4	43	44.2	4	46	35.5	4	48	32.6	4.81		38
23	13 Aug 06	B7.5	4	43	6.9	4	45	17.5	4	47	29.8	4.38		104
24	17 Aug 06	B1.0	4	28	16.8	4	29	54.9	4	32	56.8	4.67		179
25	21 Aug 06	B6.7	20	17	54.3	20	19	3	20	21	24.2	3.50		73
26	27 Aug 06	B3.3	5	37	29.5	5	38	39.1	5	42	29.4	5.00		36
27	31 Aug 06	B7.1	4	25	28.4	4	26	58.8	4	31	47.6	6.32		91

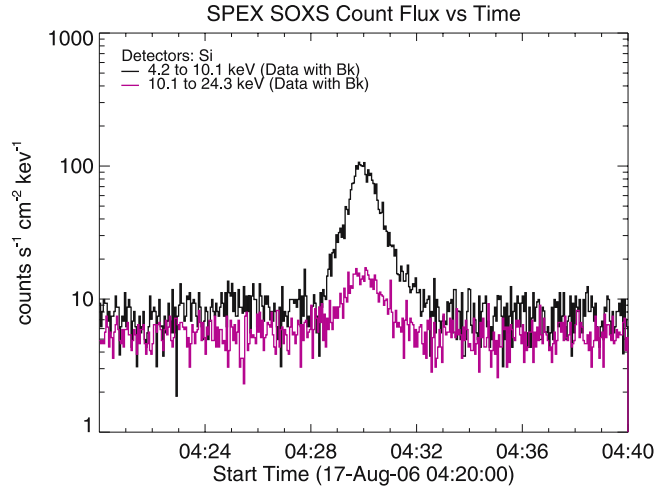


Figure 1. Light curves of the microflare observed on 17 August 2006 by Si detector of SOXS mission in 4.2–10.1 and 10.1–24.3 keV bands.

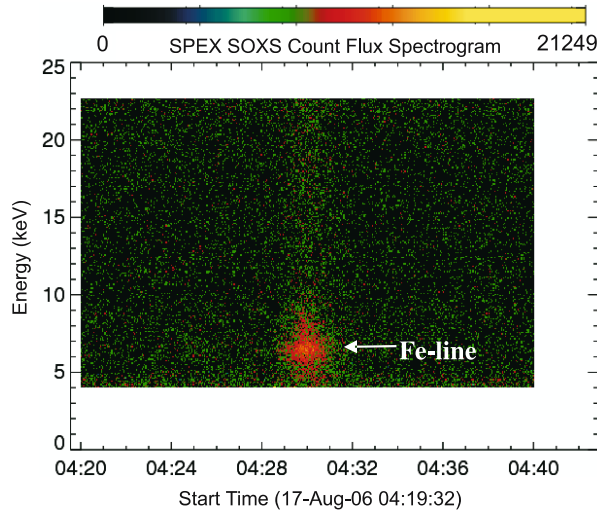


Figure 2. The spectrogram of 17 August 2006 microflare observed by Si detector of the SOXS mission. Note the Fe-line complex (red colour) shown by arrow peaking near 6.7 keV.

curves of a microflare observed on 17 August 2006 by Si detector of the SOXS in 4.2–24.3 keV. The light curves provide us the evolution and duration of the microflares. Our study of 27 microflares considered for this investigation shows the duration varies between ~ 4 and 25 min. In $H\alpha$, the microflares are classified as subflares (SF). In Fig. 2, the spectrogram of the same microflare (cf. Fig. 1) is presented, which shows unambiguously the Fe line feature indicated by arrow. It may be noted from the Fig. 2, that the microflare reveals Fe line complex (red colour) peaking at 6.7 keV. However, Fe/Ni line complex was not visible in all microflares under investigation except in two

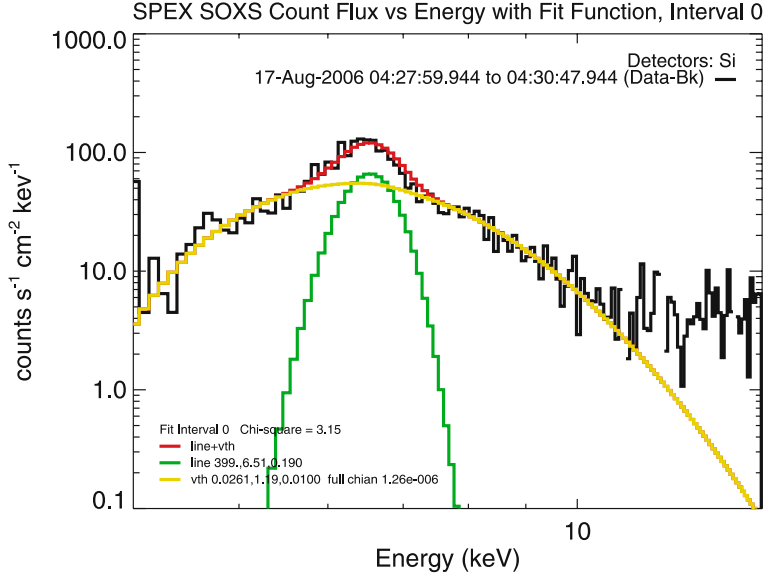


Figure 3. Count spectra of the microflare observed on 17 August 2006 by Si detector of SOXS mission and fitted with thermal and line emission.

microflares, viz., 25 February (B9.8) and 30 October 2004 (B9.2). The temperature for formation of Fe/Ni line feature has been proposed to be 15 MK (Phillips 2004; Jain *et al.* 2006a). However, the presence of Fe line feature in each microflare under study implies that the temperature in the coronal plasma had reached ≥ 9 MK, which is enough to heat the corona.

4. Discussion and conclusion

In Fig. 3, we show the count spectra of a microflare, which also shows unambiguously the presence of Fe line feature at ~ 6.7 keV. The microflares are assumed to be the result of interaction of low lying loops (Jain *et al.* 2006b; Aschwanden 2007), and thereby causing reconnection in the region ranging from the upper chromosphere to transition region or corona. The electrons accelerated during the reconnection heat the ambient plasma in the upper chromosphere, transition region or corona. They also interact with ambient material while moving down along the loops and loosing their energy to the target material to heat it during each collision. The heated plasma move up (upflows) from chromosphere or transition region to producing soft X-ray emission as revealed in the X-ray photon spectrum between 4 and 8 keV of Fig. 3 including Fe-line complex. The other possibility of microflares occurrence through reconnection is in the slightly higher altitude ($\sim 1\text{--}3 \times 10^4$ km) coronal loops and the accelerated electrons could heat the plasma at the intersection region of the separatrixes of the active region, which however, may reveal continuum X-ray emission as seen in the spectrum beyond 8 keV in Fig. 3. We analysed 2–5 spectra depending on the duration of each microflare under study and found that the best fit to the spectra is by thermal for energy range 4–8 keV, while the continuum of the spectra at energies > 8 keV fits well by broken power-law (Jain *et al.* 2008). This result shows that the microflares

are most likely the product of both mechanism viz., thermal and non-thermal, which is an indicator of hybrid mechanism in microflares as proposed earlier by Jain *et al.* (2006b). Nevertheless, our investigation enabled us to conclude that the microflares produce enough energy/temperature irrespective to their origin in upper chromosphere, transition region or corona to heat the active region coronal plasma as inferred from the presence of Fe-line feature and non-thermal spectra. The Fe has high atomic number (Z) and therefore settles down against gravitation as soon as the temperature of the active region corona goes below 9 MK. However, as inferred from our observations, during the occurrence of microflare the temperature in the active region corona exceeds 9 MK indicating that microflares are sufficient to heat the corona. It might also be possible to observe the presence of Fe-line feature in nanoflares if the detectors are designed of appropriate sensitivity.

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Initiation and Propagation of Coronal Mass Ejections

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Abstract. This paper reviews recent progress in the research on the initiation and propagation of CMEs. In the initiation part, several trigger mechanisms are discussed; in the propagation part, the observations and modelings of EIT waves/dimmings, as the EUV counterparts of CMEs, are described.

Key words. Coronal mass ejections—magnetic field.

1. Introduction

Coronal mass ejections (CMEs) have been observed for over 30 years. They keep being an intriguing research topic, not only because they are now realized to be the major driver for space weather disturbances, which are intimately connected to human activities, but also because they themselves are full of questions that have been provoking scientists to seek for answers. Stimulated by the limited observations, theoretical researches are involved in the various phases of the eruptions from their birth to their pilgrimage in the interplanetary (IP) space. First, we are still not quite sure what is the progenitor of a CME. The rough picture is described as follows: magnetic field, which is generated at the tachocline layer, emerges throughout the convection zone and the lower atmosphere into the tenuous corona. The coronal field keeps adjusting to a more and more complex magnetic structure in a quasi-steady way. After a threshold, the magnetic structure cannot sustain its equilibrium and begins to erupt. In this picture, it is still an open question whether the pre-CME structure should always possess a flux rope. Or, the so-called flux rope is actually an extreme case of the ordinary magnetic arcade with a strong twist. The second issue is how the progenitor is triggered to deviate from the equilibrium state. In this aspect, the statistical investigations of the correlation between CME onsets and other phenomena are of extreme significance. The third issue is how a CME is accelerated. The related questions involve (1) whether magnetic reconnection is a necessary condition, (2) how important the interaction between the ejecta and the solar wind is, (3) the effect of prominence mass drainage, among others. The fourth issue is how the CME is related to the accompanied phenomena, such as solar flares, Moreton waves, EIT waves and dimmings, transient coronal holes, etc. The fifth issue is how the CME evolves to an interplanetary CME (ICME) and how the CME properties affect the geomagnetic activity.

In this review paper, we focus on two aspects of the theoretical researches on CMEs, i.e., the initiation and propagation, which are presented in sections 2 and 3, respectively. We refer the readers to Forbes (2000), Gopalswamy (2003), and Volume 123 of *Space*

Sci. Rev. for more detailed reviews. The chances and challenges of solar cycle 24 are briefly prospected in section 4.

2. Initiation of CMEs

Except some narrow CMEs, which may correspond to a jet (say, reconnection jet) propagating along open field lines, most CMEs are regarded as an erupting flux rope system, with a typical three-component structure in the white-light coronagraph images, although sometimes one or two components are absent possibly due to observational effect or the plasma has not yet condensed to form a filament at the magnetic dips of the flux rope. The eruption process can generally be described in the classical CSHKP framework: a flux rope, which may or may not host a filament, becomes unstable or loses its equilibrium, it then rises and pulls up the closed field lines straddling over it, so as to form a current sheet beneath the flux rope. The reconnection at the current sheet removes the constraint of the line-tied field lines, and the flux rope is pushed to erupt by the upward reconnection jet. Therefore, one important and unclear issue in this picture is how the flux system is triggered.

2.1 Emerging flux trigger mechanism

Early in the 1970s, it was found that weak X-ray activities often precede solar flares (Datlowe *et al.* 1974), which were described as the soft X-ray precursor of CMEs by Harrison *et al.* (1985). In an apparently unrelated research, Feynman & Martin (1995) found that many CMEs are strongly associated with emerging flux that possesses polarity orientation favorable for magnetic reconnection between the emerging flux and the pre-existing coronal field either inside or outside the filament channel. Wang & Sheeley (1999) confirmed the strong correlation between CMEs and reconnection-favorable emerging flux, although it is noted that not all CMEs are related to emerging flux. Motivated by such a correlation, we proposed an emerging flux trigger mechanism for CMEs (Chen & Shibata 2000), as illustrated by Fig. 1. When the reconnection-favorable emerging flux appears inside the filament channel, it cancels the small magnetic loops near the polarity inversion line (PIL). Thereby, the magnetic pressure

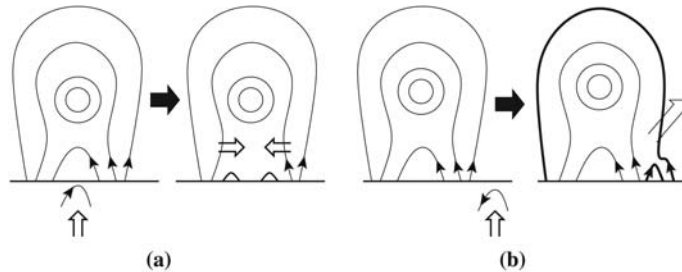


Figure 1. Schematic diagram of the emerging flux trigger mechanism for CMEs. **(a)** Emerging flux inside the filament channel cancels the pre-existing loops, which results in the *in situ* decrease of the magnetic pressure. Lateral magnetized plasmas are driven convergently to form a current sheet; **(b)** Emerging flux outside the filament channel reconnects with the large coronal loop, which results in the expansion of the loop. The underlying flux rope then rises and a current sheet forms near the magnetic null point.

decreases locally. Plasmas on both sides of the PIL, which are initially in equilibrium, are driven to move convergently along with the frozen-in anti-parallel magnetic field under the pressure gradient. As a result, a current sheet forms above the PIL, and the flux rope is also triggered to move slightly upward. The ensuing reconnection at the current sheet leads to the formation of a cusp-shaped two-ribbon flare and the fast eruption of the flux rope system. When the reconnection-favorable emerging flux appears outside the filament channel (say, on the right side), it reconnects with the large-scale magnetic arcades that cover the flux rope. The right leg of the arcade, which is rooted very close to the PIL, is re-connected far from the PIL on the right side of the emerging flux. The magnetic tension force along the curved field line pulls the arcades to move upward, with the flux rope following immediately. The rising flux rope pulls the overlying field lines up and a current sheet forms near the null point below the flux rope. Similarly, the magnetic reconnection at the current sheet leads to a two-ribbon flare and the fast eruption of a CME.

In this model, the onset of the CME is triggered by the localized reconnection between emerging flux and the pre-existing coronal field. Such a reconnection produces X-ray jets (and $H\alpha$ surges if chromosphere is considered as done in Yokoyama & Shibata 1995), which correspond to the soft X-ray precursor well before the main flare as mentioned by Harrison *et al.* (1985). The numerical results also show that the impulsive phase of the main flare coincides with the acceleration phase of the CME, and after the flare peak, the CME moves with an almost constant velocity. Jing *et al.* (2004) found that about 68% of disk CMEs are associated with emerging flux. So, the onset of quite a large part of the CMEs can be explained by our model. The simulation results in this model were also found to be consistent with various observations (e.g., Zhang *et al.* 2001). In particular, Sterling & Moore (2005) analyzed a CME event, in which they found that the height profile of the filament is very similar to that in our paper. A parameter survey of this model was conducted by Xu *et al.* (2005); its image synthesis was composed by Shiota *et al.* (2005) in order to compare with Yohkoh/SXT images. Such a model was recently extended to the spherical co-ordinators (Dubey *et al.* 2006).

2.2 Other trigger mechanisms

Observations have shown various kinds of evolving magnetic structures, for example, converging motion of the filament channel (note that the apparent converging motion may be a result of the diffusion of magnetic polarities), shear motion, twist motion, decay of the active region, etc. Accordingly, trigger mechanisms based on these changes have also been proposed. van Ballegoijen & Martens (1989) proposed that the converging motion of magnetic arcades, by which a filament may be formed, can also lead to the destabilization of the filament; Mikić *et al.* (1988) found that after a large enough shear, the closed magnetic arcades would asymptotically approach the open field, while a resistive instability can result in the eruption. Kusano *et al.* (2004), however, found that reversed magnetic shear could also trigger the eruption. Both analytical and numerical simulations indicate that there may exist catastrophic behavior in the flux rope motions as the footpoints of the magnetic arcades converge or shear (Forbes & Priest 1995; Hu & Jiang 2001). The analytical investigation by Isenberg *et al.* (1993) illustrates that the gradual decay of the background magnetic field would also cause the flux rope to lose equilibrium catastrophically. In all of these

three cases and our emerging flux model, the essence is that the evolving magnetic structure either increases the magnetic pressure below the flux rope or decreases the magnetic tension force above the flux rope, thereby the flux rope cannot sustain its equilibrium.

Chen *et al.* (1997) and Krall *et al.* (2001) proposed that the injection of poloidal magnetic flux into the flux rope would cause the flux rope to erupt. Physically this process is similar to the kink instability model as put forward by Hood & Priest (1981) and simulated by Török and Kliem (2005). The model has been compared with observations in many cases. A modified version of the kink model, i.e., the rupture mechanism, was proposed by Sturrock *et al.* (2001) and simulated by Fan (2005), where part of the flux rope penetrates the overlying magnetic field and erupts into the IP space.

In order to circumvent the Aly's constraint, Antiochos *et al.* (1999) proposed a magnetic breakout model, i.e., only the sheared part of the closed field lines near the PIL is opened during the CME. The essence of this model is that the overlying background magnetic field reconnects with the sheared arcade at the magnetic null point above the latter, by which the constraint over the sheared arcade is removed gradually like an onion-peeling process. If such a reconnection above the arcade exists during the onset of the CME, it is expected to see soft X-ray bright loops on both sides of the sheared arcade and reverse type III radio bursts that are produced by the reconnection-accelerated electrons.

There are some other less-recognized trigger mechanisms for CMEs. Filament mass drainage, by which the filament obtains a buoyancy force, may play a role in triggering the onset of a CME (Low 2001), which was identified in one event recently (Zhou *et al.* 2006). Moreton and EIT waves, which are generated by a remote CME, often trigger a filament to oscillate, and erupt sometimes (Ballester 2006), which deserves further investigations.

3. Propagation of CMEs

As mentioned in section 1, there are many interesting topics related to the propagation of CMEs. Here, we just mention the EUV counterparts of propagating CMEs, i.e., EIT waves/dimmings.

3.1 EIT wave/dimming observations

EIT waves were originally observed by the EIT telescope onboard the SOHO satellite as propagating wave-like fronts, with an emission enhancement ranging from 25% to less than 14%, which is followed immediately by expanding EIT dimmings (Thompson *et al.* 1998). Therefore, EIT waves and dimmings are symbiotic phenomena. One typical feature is that the bright fronts propagate only in the quiet regions, avoiding any active region. Therefore, when the large-scale magnetic configuration is simple, for instance, with only one active region on the visible disk, the EIT wave fronts are almost circular; however, when there are other active regions surrounding the source region of the eruption, the EIT waves appear in patches, managing their ways outward separately in the quiet regions.

Wills-Davey & Thompson (1999) found that EIT waves can be observed in both 195 Å (with the formation temperature $T \sim 1.4$ MK) and 171 Å (with the formation

temperature $T \sim 1$ MK), with more detailed structures in the $171 \text{ \AA}$ images. Later, Zhukov & Auchere (2004) also identified EIT waves in $284 \text{ \AA}$ (with the formation temperature $T \sim 1.9$ MK). Since the brightenings are observed at very different temperatures, it is concluded that they are mainly due to the density enhancement, although Wills-Davey & Thompson (1999) and Chen & Fang (2005) pointed out that temperature effect is not negligible. Weak dimmings are also reported at $304 \text{ \AA}$ (Chertok & Grechnev 2003). However, it is not sufficient to say they have imprints in the chromosphere since a coronal line Si XI $303.32 \text{ \AA}$ and a transition region line He II $303.78 \text{ \AA}$ are blended at the EIT 304 bandpass. It is generally believed that EIT waves are a phenomenon propagating in the corona. Based on the observational results in Thompson *et al.* (2000) and Harrison *et al.* (2003), Chen & Fang (2005) proposed that the “EIT waves” map the footprints of the CME leading edge, and the dimming region maps the bottom of the CME cavity. Therefore, EIT waves/dimmings are actually the EUV counterparts of CMEs.

3.2 Debates on EIT wave mechanism

Early in the 1960s, it was discovered in the $H\alpha$ line wing that arc-shaped chromospheric perturbations propagate away from some big flares (Moreton & Ramsey 1960), which were later called Moreton waves. Such a wave, with a surprisingly large velocity of the order of 1000 km s^{-1} , was later explained by Uchida (1968) as a fast-mode wave in the corona, which sweeps the chromosphere as it propagates. Since then, it was expected to detect such a wave in the corona. It was not successful except one event observed by OSO 7 satellite (Neupert 1989). The discovery of EIT waves by SOHO/EIT, then, sparked a lot of interest, as well as controversies. It was very natural to consider the EIT waves as the ever-missing coronal counterparts of Moreton waves (or coronal Moreton waves for short), i.e., they are coronal fast-mode waves. After extrapolating the coronal magnetic field based on a potential field model, Wang (2000) and Wu *et al.* (2001) claimed that the propagating fast-mode waves in the corona can match the observed EIT wave fronts. However, it is very difficult for the fast-mode wave model to explain the typical features of EIT waves: (1) The EIT wave speeds are 3 or more times smaller than Moreton waves (Klassen *et al.* 2000); (2) Delannée & Aulanier (1999) found that EIT waves stop at the magnetic separatrix, which led them to speculate that EIT waves could be associated with magnetic rearrangement; (3) The EIT velocities are not correlated with the speeds of the type II radio bursts, the latter of which are believed to be the radio signature of the coronal fast-mode shock waves; (4) EIT wave speeds can be as low as 50 km s^{-1} (Thompson & Myers 2007), which is even below the sound speed in the corona. However, fast-mode wave speed should always be larger than the sound speed.

In order to reconcile all these discrepancies, Chen *et al.* (2002 and 2005) predict that there should exist two EUV waves associated with a CME event, i.e., the coronal Moreton wave and the EIT wave, which was later confirmed by Harra & Sterling (2003). In our model, the coronal Moreton waves correspond to the piston-driven shock over the CME rather than the blast wave from the pressure pulse in the flare, and EIT waves are generated by successive opening (or stretching) of closed field lines, which is pushed by the erupting flux rope. Each field line is pushed to expand at its top, and the deformation is transferred down to the footpoints of the field line. Whenever the leg of a field line expands, the plasma outside the field line is compressed to form

an EIT wave front, while the plasma inside is evacuated, resulting in EIT dimmings. Therefore, the model can explain both EIT waves and dimmings. The numerical results reproduce many characteristics that are obtained from observations: (1) EIT waves propagate with a velocity ~ 3 times smaller than the coronal Moreton waves; (2) EIT waves stop at the magnetic separatrix between the source active region and another active region; (3) the EIT wave speed is anti-correlated with the speed of type II radio bursts.

3.3 Significance of EIT wave/dimming observations

- EIT waves/dimmings are the disk signatures of CMEs: Biesecker *et al.* (2002) found that whenever there is an EIT wave, there should be a CME in the coronagraph images, although the contrary is not true. As mentioned above, EIT waves/dimmings map the CME leading edge/cavity, and they are the disk signatures of the CMEs. Therefore, routine observations of EIT waves/dimmings will be an efficient way to monitor CMEs, especially those directed towards our earth.
- EIT dimmings provide an estimate of the mass supply for CMEs: CMEs are the major drivers for space weather disturbances such as geomagnetic storms. Their mass, as well as their velocity and the magnetic field, is an important factor that may influence their geomagnetic effect. Hence, the estimate of their mass in the early phase of the eruption is crucial for space weather forecast. Harrison & Lyons (2000) proposed that EIT dimmings, which are due to the plasma evacuation as in our model, can be used to estimate the CME mass.
- Large-scale coronal magnetic field can be inferred: Various efforts have been put into the measurement of the coronal magnetic field, such as the radio diagnosis, near infrared Zeeman effect measurements, and so on. Before these methods become practical, EIT waves/dimmings can provide an efficient way to diagnose the coronal magnetic field. As discussed in Chen *et al.* (2002 and 2005), EIT waves/dimmings are produced by the opening of the closed field lines covering the erupting flux rope. This means that coronal field lines should be self-closed within the dimming regions. With sufficiently high cadence of the EIT wave observations, their velocity pattern can even be used to derive the coronal magnetic field.

4. Prospects for the Solar Cycle 24

It is seen from the above review that our understanding of the CME initiation and propagation strongly relies on observations. For the CME initiation, it will be vital to detect the progenitor of the CME and its early evolution in order to distinguish between various trigger models. In this sense, UV coronagraph observations would be invaluable to trace the early evolution of any ongoing eruption; on the other hand, sub-surface detections of the magnetic field and motions based on the local seismology would also be helpful. For the EIT waves/dimmings, we believe that the ongoing STEREO/SECCHI observations with a high cadence would gradually uncover the veil over the spectacular phenomenon, which can then be used as a proxy for the coronal magnetic field diagnostics.

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UV Diagnostics for the Energy Budget of Flares and CMEs

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Abstract. Solar flares and coronal mass ejections convert large amounts of magnetic free energy into thermal, kinetic and potential energies, and into energy of non-thermal particles. The partitioning among these forms of energy is fundamental to both the physics of the eruptive events and the space weather consequences of the eruptions. This talk describes some aspects of the energy budget that can be derived from ultraviolet observations of the corona.

Key words. Sun: corona—flares—coronal mass ejections.

1. Introduction

The energy budget of solar flares and coronal mass ejections (CMEs) is the fundamental information needed to reach a physical understanding of the process. It is also the key information needed to apply our knowledge of solar activity to other astrophysical systems, ranging from solar-type stars to accretion disks around neutron stars and massive black holes.

It has proven very difficult to obtain a reliable energy budget for solar flare-CME events. No single wavelength band contains good diagnostics for all plasma conditions present, because gas from photospheric temperatures to tens of millions of Kelvins is present, and because non-thermal particles at very high energies contain a significant fraction of the energy. It is also very difficult to accurately measure the magnetic free energy. Finally, one must be careful not to double count components of the energy, as conversion of energy among magnetic, thermal, kinetic, radiative and non-thermal forms is the essence of the process.

The terms of the energy budgets have been compared in a few cases. Emslie *et al.* (2005) estimated the flare thermal and non-thermal energies and the CME kinetic and solar energetic particle (SEP) energies for two X-class events. They found that the magnetic, flare and CME energies were comparable, and that the flare energy in non-thermal particles was similar to that of the thermal plasma. Subramanian & Vourlidas (2007) measured the kinetic energies for several CMEs and estimated the magnetic energies from a magnetic flux rope model. They found comparable magnetic and kinetic energies. Akmal *et al.* (2001) and Ciaravella *et al.* (2001) compared the kinetic energies of CMEs with the heat that went into maintaining the plasma temperature in the face of adiabatic cooling, and they found that the two values were comparable.

Here we consider how ultraviolet spectroscopy can provide measurements of some of the terms in the energy budget. The ultraviolet wavelength range contains diagnostics for plasma as hot as 10 MK, and comparison of different spectral lines provides additional information about the line-of-sight component of velocity, the plasma composition, and through the ionization state, the thermal history of the plasma. Very few UV spectra of flares or of CMEs were available before the launch of SOHO, but now there are spectra covering a wide range of temperatures and velocities from CDS (e.g., Pike & Mason 2002), SUMER (e.g., Innes *et al.* 2001) and UVCS (see review by Kohl *et al.* 2006).

Here we discuss two applications of the UV spectroscopy to the energy budget problem; the radiative energy losses during the impulsive phase of an X-class flare, and the post-ejection heating of CME plasma as it travels away from the Sun.

2. Transition region emission in the impulsive phase of flares

During the impulsive phase of a solar flare, magnetic energy is transformed into several other forms, mainly the kinetic energy of non-thermal particles, heat of the flare plasma, electromagnetic radiation and kinetic energy of mass motions including coronal mass ejections and turbulence. Understanding the partitioning of energy among these forms is a necessary step to understand the physical processes involved. In large flares, the energy in non-thermal electrons inferred from the bremsstrahlung spectrum of hard X-rays is comparable to the thermal energy of the plasma that produces the $\sim 2 \times 10^7$ K soft X-ray emission (Holman *et al.* 2003; Emslie *et al.* 2005) and the kinetic energy of plasma ejected in the CME is comparable to those (Schwenn *et al.* 2006).

There are some difficulties in making this comparison, however. First, the thermal energy of the X-ray emitting gas is

$$E_{th} = \frac{3}{2}(n_e + n_p)kTV, \quad (1)$$

where n_e and n_p are the electron and proton densities, T is the temperature, and V is the volume. The temperature is measured from the soft X-ray spectrum, and the product $n_e n_p V$ is the emission measure, which can be derived from the luminosity and temperature. The difficulty is that there is seldom an accurate measure of the density or volume. It is sometimes assumed that V is the apparent volume estimated from X-ray images, but it is likely that the real volume is a small fraction of the apparent volume, given by a filling factor f (e.g., Aschwanden & Aschwanden 2006). If the filling factor is 0.1, the thermal energy estimated from the apparent volume is too large by a factor of 3.

The second difficulty is that a significant fraction of the thermal energy might be carried away from the X-ray emitting plasma by thermal conduction and radiated at lower temperatures. Estimates of the thermal conduction losses suffer from an uncertainty in filling factor, this time the areal filling factor of hot flux tubes at transition region temperatures, and from the possibility that classical thermal conduction as given by the Spitzer conductivity does not apply.

Ultraviolet spectra have the potential to address both of these questions, because there are density-sensitive line ratios in the UV and because most of the energy carried away by thermal conduction would be radiated away in UV lines such as Ly α and

O VI $\lambda\lambda 1032, 1038$. Because most UV spectrometers rely on an entrance slit to separate spatial from spectral information, however, it is unlikely that a spectrograph will be pointed at exactly the right place at the right time. A raster scan of the spectrograph slit can in principle give adequate spatial and spectral coverage, but in practice the raster cadence for the SOHO instruments was on the order of half an hour.

Here we address the thermal conduction problem by observing photons from the flare that are scattered off ions in the corona above the flare, following the methods and results presented by Raymond *et al.* (2007). That way, we measure the total flare luminosity during the impulsive phase in O VI and Ly α lines. Ly α is formed at about 20,000 K, and it dominates the luminosity of the upper chromosphere and lower transition region. O VI is formed at about 300,000 K, near the peak of the radiative cooling curve. The O VI doublet accounts for about 1/8 of the emission at that temperature for solar photospheric abundances, or about 1/16 the total emission between 10^5 and 10^6 K for an emission measure slope like that seen in the quiet Sun or active regions (e.g., Raymond & Doyle 1981).

Figure 1 shows the UVCS observation for the flare of 4 November, 2003, which saturated the GOES detectors, but was at least class X-17. The panel shows the brightness of the O VI $\lambda 1032$ line as a function of spatial position along the slit (horizontal direction covering position angles between 303 and 230 degrees from solar North) and time (vertical). The brightening that appears at 19:42 UT (along with a somewhat fainter brightening at 19:33 UT) results from O VI photons scattered from O VI photons in the corona. Immediately after the brightening, the O VI emission fades. This comes about because the CME reaches the position of the UVCS slit, which was placed at $1.63 R_{\odot}$ during this set of observations, and blows the pre-CME corona away.

We know that the O VI brightening results from scattered flare photons for several reasons. First, the emission above the pre-CME background shows a 4:1 intensity

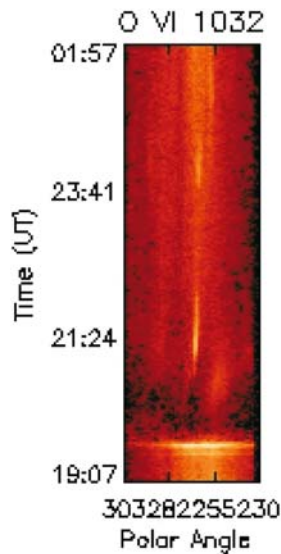


Figure 1. Brightness of the O VI $\lambda 1032$ line as a function of time (vertical axis) and position along the UVCS slit.

Table 1. November 4, 2003 energy budget (10^{26} erg s^{-1}).

Event	T_{start}	L_{TR}	L_{therm}	dE/dt
November 4, 2003	19:31:12	2.7	0.7	32
	19:33:22	18.5	7.7	155
	19:35:32	12.2	15.6	-6

ratio between the $\lambda 1032$ and $\lambda 1038$ members of the doublet. This is the value for scattered disk radiation, as opposed to a 2:1 ratio for collisional excitation. Second, the emission appears simultaneously all along the slit, while anything related to CME plasma happens first at the center of the slit and spreads along the slit over the course of ~ 5 minutes. Third, the O VI emission peaks simultaneously with the RHESSI hard X-ray emission. Fourth, the O VI line profile widths and centroids do not change from their pre-CME values.

Several steps are required to derive the luminosity from gas at transition region temperatures. We determine the column density along the line of sight for each pixel along the UVCS slit from the pre-CME intensities of the O VI lines, the scattering cross section and the intensity of O VI emission from the disk (Raymond *et al.* 2007). Then from the column density, the added flux during the flare, and the scattering cross section we compute the luminosity in O VI photons. Finally, we use model calculations of the ratio of O VI luminosity to total transition region luminosity to find L_{TR} . There are a few interesting technical details, most notably the effects of a Doppler shift of the flare emission. Doppler shifts of 20 to 100 km s^{-1} are reported in flares (e.g., Czaykowska *et al.* 1999; Milligan *et al.* 2006). This shifts the flare photons partly away from the absorption profile of the scattering ions in the corona. The effects of the Doppler shift can be determined by comparing the scattered intensity at different places along the UVCS slit, since only the component of velocity along the direction between the flare and the scattering plasma affects the scattering. Among the flares analyzed so far, the Doppler shift ranges up to 90 km s^{-1} .

The overall result of this analysis is that the $Ly\alpha$ and transition region luminosities are comparable. They are on the order of 10% of the rate of increase of the thermal energy of the X-ray emitting plasma if the filling factor of the hot gas is near 1, but they could be comparable if the filling factor is on the order of 0.01.

Table 1 shows the power in different terms in the energy budget for the November 4, 2003 flare in two-minute intervals corresponding to the UVCS exposures. L_{TR} is the transition region luminosity derived from the UVCS observation, L_{therm} is the luminosity of the X-ray emitting thermal plasma, and dE/dt is the rate of change of the thermal energy of the X-ray emitting plasma assuming a filling factor f of 1.0 (Raymond *et al.* 2007). The table shows that the transition region luminosity exceeds the X-ray luminosity for a few minutes during the impulsive phase, but that the X-ray luminosity comes to dominate as cool material is evaporated into the corona.

3. Post-ejection heating

The mass and kinetic energy of CME plasma can be determined from a series of white light coronagraph images (Subramanian & Vourlidas 2007), but the thermal energy is more difficult to determine. UV spectra provide upper limits to the proton and ion

temperatures through the measured line widths, and they provide electron temperatures because of the temperature dependence of the excitation rates of the lines. Moreover, the ionization state of the plasma freezes in, often before the plasma reaches the UVCS slit, so that a UV spectrum gives information about the thermal history of the plasma.

In general, the plasma that shows up brightly in UVCS spectra is relatively cool, typically 50,000 to 500,000 K. That means that its thermal energy is relatively low, $\sim 2 \times 10^{13}$ erg/g, compared to the kinetic energy of 5×10^{15} erg/g for a 1000 km s⁻¹ CME. However, the gas observed by UVCS has expanded by many orders of magnitude by the time it reaches the UVCS slit, and adiabatic cooling would make it much cooler than the observed temperature range if the gas were not heated long after it is ejected from the solar surface. By the time it reaches 1 AU, the temperature is still on the order of 10^5 K, so considerable ongoing heating is required to maintain the temperature.

There are few theories available for the heating of CMEs after the eruption, or for that matter for heating of the plasma during the eruption itself. UVCS is most sensitive to dense, relatively cool material that emits very strongly in ions such as C III, which are completely absent at coronal temperatures. Therefore, ejected prominence material that has not been too strongly heated is the easiest to detect. However, hotter gas is observed by means of O VI, Si XII (2×10^6 K) and [Fe XVIII] ($5\text{--}6 \times 10^6$ K). The last of these represents plasma heated by reconnection during the flare itself, while the O VI and Si XII could be ejected active region material.

One plausible source of heat is the CME-driven shock wave (Raymond *et al.* 2000; Mancuso *et al.* 2002), but that ultimately derives from the CME kinetic energy, and it cannot reach the cool components considered here, so we will not consider it for the energy budget. Another is thermal conduction from hotter coronal material. This may be important for heating electrons at solar wind distances beyond 0.1 AU, since the CME plasma is more or less like relatively fast solar wind at that point. It has more difficulty explaining heating in plasma below 10^5 K, though, because of the strong temperature dependence of the thermal conductivity. Another possibility is wave heating similar to that that heats streamers and drives the quiescent solar wind.

The most interesting possibility was proposed by Kumar & Rust (1996), who suggested that conservation of magnetic helicity in an expanding CME implies the dissipation of magnetic energy. They found that expansion of a self-similar magnetic flux rope transforms comparable amounts of magnetic energy into kinetic energy and plasma heating. They did not really identify the heating mechanism, but they mentioned a turbulent cascade as a likely candidate. More recently, Liu *et al.* (2006) determined the heating rates for interplanetary CMEs at distances between 0.3 AU and 20 AU. They found that heating of protons and alpha particles is consistent with the dissipation of turbulence, and they showed that temperature anisotropies do not drive the turbulence in the inner heliosphere, but they were not able to pin down the source of turbulent energy.

Heating rates have been determined with UVCS in a few cases (Akmal *et al.* 2001; Ciaravella *et al.* 2001; Lee *et al.* 2007). The method is to start with the density either from a density sensitive line ratio such as the ratio of the [O V] forbidden line at 1213.8 Å to the O V] intercombination line at 1218.4 Å, or else from the electron column density from LASCO divided by the line of sight depth from a 3D reconstruction. The temperature and ionization state of the plasma are constrained by the intensities of emission lines such as Ly α , C III and O VI. A large grid of models with different

Table 2. CME energy budgets (10^{13} erg/gm).

	C III ^a	O V ^a	O VI ^a	AH&H ^b	K&R ^b
Kinetic	180	180	180	160	160
Gravitational	78	78	78	56	56
Ionization	1.9	2.1	2.1	2.1	2.1
Thermal	1–3	2–5	3–10	3–4	3.4
Total heat	10–50	50–200	10–500	920–1400	570–690

^aApril 23, 1999 CME.^bDecember 13, 2001 CME.

initial temperatures and densities and with different parameterized heating rates is computed, and for a CME feature with good enough observations, nearly all those models can be eliminated. Note that any heating of the plasma during the very early stages of the eruption is included in the initial temperature specified. Such heating tends to strongly affect the ionization state of the plasma, but it is rapidly lost to adiabatic expansion.

The integrated heating of the models that match the observations tends to lie in a range of about an order of magnitude for the Akmal *et al.* (2001) and Ciaravella *et al.* (2001) analyses or somewhat less for the cases analyzed by Lee *et al.* (2007). The total heat that must be added to the plasma after ejection is comparable to or larger than the kinetic energy as shown in Table 2. The results for the April 23, 1999 and December 13, 2001 events are taken from Akmal *et al.* (2001) and Lee *et al.* (2007), respectively. The first three columns show results for 3 different knots of plasma bright in C III, O V and O VI, respectively, for the April 1999 event at $3.5 R_{\odot}$ (Akmal *et al.* 2001). The last two columns show results for one knot for the December 2001 event at $2.9 R_{\odot}$ for heating rates given by the functional forms of Allen *et al.* (1998) and of Kumar & Rust (1996) indicated by AH&H and by K&R, respectively (Lee *et al.* 2007).

Overall, we find that the heat input to CME plasma after ejection from the Sun is comparable to the kinetic energy in all but the coolest knots, and larger than the kinetic energy in some cases. This energy is important for a complete physical understanding of CMEs, but we must still narrow down the possible sources of that energy.

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X-ray Emission Characteristics of Flares Associated with CMEs

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Abstract. We present the study of 20 solar flares observed by “Solar X-ray Spectrometer (SOXS)” mission during November 2003 to December 2006 and found associated with coronal mass ejections (CMEs) seen by LASCO/SOHO mission. In this investigation, X-ray emission characteristics of solar flares and their relationship with the dynamics of CMEs have been presented. We found that the fast moving CMEs, i.e., positive acceleration are better associated with short rise time (< 150 s) flares. However, the velocity of CMEs increases as a function of duration of the flares in both 4.1–10 and 10–20 keV bands. This indicates that the possibility of association of CMEs with larger speeds exists with long duration flare events. We observed that CMEs decelerate with increasing rise time, decay time and duration of the associated X-ray flares. A total 10 out of 20 CMEs under current investigation showed positive acceleration, and 5 of them whose speed did not exceed 589 km/s were associated with short rise time (< 150 s) and short duration (< 1300 s) flares. The other 5 CMEs were associated with long duration or large rise time flare events. The unusual feature of all these positive accelerating CMEs was their low linear speed ranging between 176 and 775 km/s. We do not find any significant correlation between X-ray peak intensity of the flares with linear speed as well as acceleration of the associated CMEs. Based on the onset time of flares and associated CMEs within the observing cadence of CMEs by LASCO, we found that in 16 cases CME preceded the flare by 23 to 1786 s, while in 4 cases flare occurred before the CME by 47 to 685 s. We argue that both events are closely associated with each other and are integral parts of one energy release system.

Key words. X-ray flares: rise time, duration, peak intensity—CME: linear velocity, acceleration.

1. Introduction

Coronal mass ejections (CMEs) have been largely known for their impact on the earth’s environment. The complex relationship between solar flares and CMEs has

been studied extensively to understand the cause and consequence between them but so far no satisfactory result has been obtained. Although it has been known that the two phenomena often occur in conjunction, the relationship is not one-to-one and the exact nature of the flare and CME triggers is not known. Whilst statistical studies indicate that higher intensity events are more likely to be accompanied by a CME (Harrison 1995), this is not always the case. For example, Green *et al.* (2002) present analysis of an X1.2 class flare for which there was no associated CME.

Kahler (1992) pointed out if the CME is associated with flare then the CME originates in the explosive phase of the flare and such flares are long-decay events (LDEs); but the relationship of CMEs with impulsive flares is still unknown. Gosling (1993) showed that solar flares play no fundamental role in causing geomagnetic disturbances. He cited studies which indicate that CMEs are the primary cause of these disturbances. The probability of CME–flare association increases with flare duration (Sheeley *et al.* 1983): 26% for duration < 1 h and 100% for duration > 6 h. It must be pointed out that some major flares associated with large-scale CMEs are not long-duration events (Nitta & Hudson 2001; Chertok *et al.* 2004). Currently, there are three ideas about the flare–CME relationship:

1. Flares produce CMEs (Dryer 1996).
2. Flares are a by-product of CMEs (Hundhausen 1999), and
3. Flares and CMEs are part of the same magnetic eruption process (Harrison 1995; Zhang *et al.* 2001).

Studies on temporal correspondence between CMEs and flares have concluded that CME onset typically precedes the associated X-ray flare onset by several minutes (e.g., Harrison 1991). This observational fact is considered to be a serious difficulty for flares to produce CMEs (Hundhausen 1999). The reconnection that leads the flare and also forms post-flare loops can be thought of as the force that propels overlying loops as CMEs (Anzer & Pnueman 1982). Kahler *et al.* (1989) argued against such a model because they could not find evidence for a flare impulsive phase affecting the height-time history of CMEs. Zhang *et al.* (2001) investigated four CMEs and compared their time evolution with GOES X-ray flares. They found that the CMEs started accelerating impulsively until the peak of the soft X-ray flare, consistent with an earlier result that flare-associated CMEs are in general faster than other CMEs (MacQueen & Fisher 1983).

According to Harrison (1995), the flare and CME are signatures of the same magnetic complexity, which represent the responses in different parts of the magnetic structure, to a particular activity, though they do not drive one another but are closely related. Su *et al.* (2007) examined the correlations between the soft X-ray peak flare flux (PFF), the CME speed (V_{CME}) and six magnetic parameters of the flaring active region. These six parameters were: the average background magnetic field strength (B), the area of the region where B is counted (S), the magnetic flux of that region (ϕ), the initial shear angle (θ_1 , measured at the flare onset), the final shear angle (θ_2 , measured at the time when the shear change stops), and the change of shear angle ($\theta_{12} = \theta_1 - \theta_2$) of the footpoints. They found no correlation between θ_1 and the intensity of flare/CME events, while the other five parameters were either positively or negatively correlated with both $\log_{10}$ (PFF) and V_{CME} . Among these five parameters, ϕ and θ_{12} showed the most significant correlations with $\log_{10}$ (PFF) and V_{CME} . The fact that both $\log_{10}$ (PFF) and V_{CME} are highly correlated with θ_{12} rather than θ_1 indicates that the intensity

of flare/CME may depend on the released magnetic free energy rather than the total free energy stored prior to the flare. Torok & Kleim (2007) concluded that increased magnetic complexity, reflected on steep magnetic gradients in the source AR's corona, tends to produce faster CMEs. Good correlations have been found between different parameters representing the magnetic shear (or twist) or the non-potentiality of the active region and the flare/CME productivity (Falconer *et al.* 2006; Jing *et al.* 2006, and references therein).

This study aims to investigate the temporal correspondence of the X-ray emission characteristics of the flares with the dynamics of the associated CMEs. The unambiguous identification of X-ray emission characteristics with dynamics of CMEs will improve the understanding of the physical processes that link the solar and interplanetary relationship.

2. Data analysis and results

We have chosen the X-ray flares observed by Solar X-ray Spectrometer (SOXS) experiment onboard GSAT-2 Indian spacecraft (Jain *et al.* 2005, 2006) during November 2003 to December 2006 and those found unambiguously associated with coronal mass ejection (CME) observed by LASCO/SOHO. Our daily X-ray observations and the preliminary light curves revealed by Si and CZT detectors in the energy range 4–25 and 4–55 keV respectively are presented at the SOXS website (<http://www.prl.res.in/~soxs-data/>). The observations of CMEs and preliminary kinematics are presented in the LASCO/CME Catalog at the website (http://cdaw.gsfc.nasa.gov/CME_list/). This catalog contains all the CMEs detected by the LASCO coronagraphs C2 and C3, which cover a combined field of view of 2.1 to 32 Rs. The actual increase of CME's height (in units of Rs) with time as it expands away from the Sun is available in the text file on the website. Each CME is characterized by the linear speed obtained by fitting a straight line (a linear or first-order polynomial fit) to the height-time measurements. The second-order fit gives the average acceleration of the CME. The acceleration of a CME can be positive, negative or close to zero suggesting CMEs speed up, move with constant speed or slow down within the FOV of the LASCO instrument. A minimum of three height-time measurements are required for an estimate of the acceleration, but the accuracy increases with more measurements. We succeeded in choosing 20 solar flare events clearly associated with CMEs for the current investigation and their temporal characteristics are presented in Table 1. The flares have been observed by both Si and CZT detectors of SOXS experiment in 4–25 and 4–55 keV respectively. The temporal characteristics and the peak flux of the flare observed in 4.1–10 keV of the CZT detector are presented in Table 1. The magnitudes of the flares range between B2.0 and M6.4 of GOES intensity class. The CME central position angle (CPA), angular width, the plane-of-sky velocity and acceleration magnitude are also shown in Table 1.

Since hard X-ray emission is one of the most significant properties of solar flares we exploit this emission in current investigation to better understand the relation between flares and CMEs in addition to soft X-ray emission in 4.1 to 10 keV band. Figure 1 shows the evolution of the solar flares observed by CZT detector of SOXS on 30 July 2003, 31 October 2004 and 25 August 2005 respectively in the energy bands 4.1–10, 10–20, 20–30 and 30–55 keV as a function of time.

Table 1. Summary of solar flares and associated CME observations.

Date	Flare characteristics				CME characteristics					
	Time			GOES class	Peak flux*	Onset time	CPA (Deg.)	Angular width (Deg.)	Linear speed (km/s)	Acceleration (m/s ²)
	Onset	Peak	End							
30 July 03	4:07:58	4:09:39	4:29:18	M2.5	3842	4:06:40	55	9	589	58.7
13 November 03	5:00:23	5:01:48	5:15:49	M1.6	1182	5:00:00	105	62	598	-5.9
7 January 04	3:55:41	4:01:12	4:45:00	M4.5	3932	3:35:50	78	171	1581	-60.4
8 January 04	4:56:36	5:03:00	5:34:42	M1.3	371	4:45:20	83	144	1713	-14.6
19 March 04	4:13:19	4:15:36	4:24:39	C1.2	22.3	4:00:00	52	37	302	31.2
5 April 04	5:37:41	5:50:57	6:35:29	M1.7	1607	5:19:48	111	191	608	-9.7
11 April 04	3:57:13	4:08:44	4:56:25	C9.6	103	3:58:00	203	314	1645	-77.6
16 June 04	4:28:16	4:33:03	4:58:01	C2.8	50	4:10:00	72	127	603	-2
26 June 04	4:17:11	4:22:48	4:30:54	C1.3	29.2	4:20:50	250	69	775	7.5
21 July 04	4:55:20	4:57:47	5:02:37	C8.9	23.3	4:47:00	165	65	419	11.4
12 August 04	4:41:38	4:43:58	4:58:30	M1.2	56.2	4:11:52	198	115	176	-2
17 August 04	5:01:05	5:02:40	5:37:26	M1.1	271	5:12:30	116	131	637	14.4
19 August 04	4:12:10	4:15:54	4:24:23	B2	20.1	3:50:00	103	39	279	6.5
23 October 04	3:58:35	4:00:05	4:05:25	C1.8	70.4	4:00:00	99	61	468	14.4
31 October 04	5:21:35	5:31:02	5:54:04	M2.3	4209	5:05:00	251	62	265	7.6
12 July 05	4:15:23	4:28:19	4:38:19	B6.0	28	4:08:20	280	34	711	-3.9
27 July 05	4:38:35	4:53:34	5:30:22	M3.7	3096	4:38:00	Halo	360	1787	-75.4
3 August 05	4:57:18	5:04:12	5:24:50	M3.4	3811	4:40:00	104	65	479	-6.7
25 August 05	4:33:05	4:39:25	5:04:11	M6.4	7346	4:25:00	115	146	1327	-66.5
3 November 05	4:33:41	4:44:13	4:53:37	B9.1	15.7	4:13:48	155	77	304	21.6

* (Counts/s/cm²/keV in 4.1–10 keV).

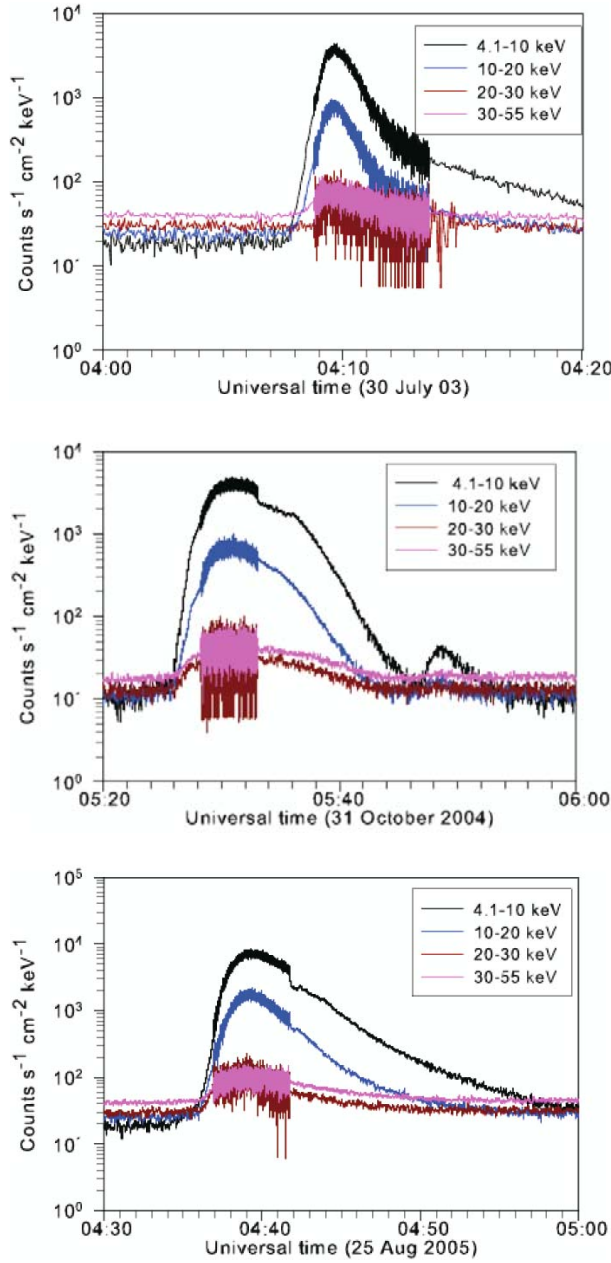


Figure 1. Light curves of 30 July 2003 (top), 31 October 2004 (middle) and 25 August 2005 (bottom) solar flares as observed by CZT detector of SOXS in 4.1–10, 10–20, 20–30 and 30–55 keV energy ranges.

In preview that the relative intensity on time scale may provide important information on the heating and cooling processes of the coronal flare plasma we measure rise time, decay time and total duration of each flare in soft and hard X-ray bands viz., 4.1–10 and 10–20 keV respectively. Start time of the flare is acquired as the time when

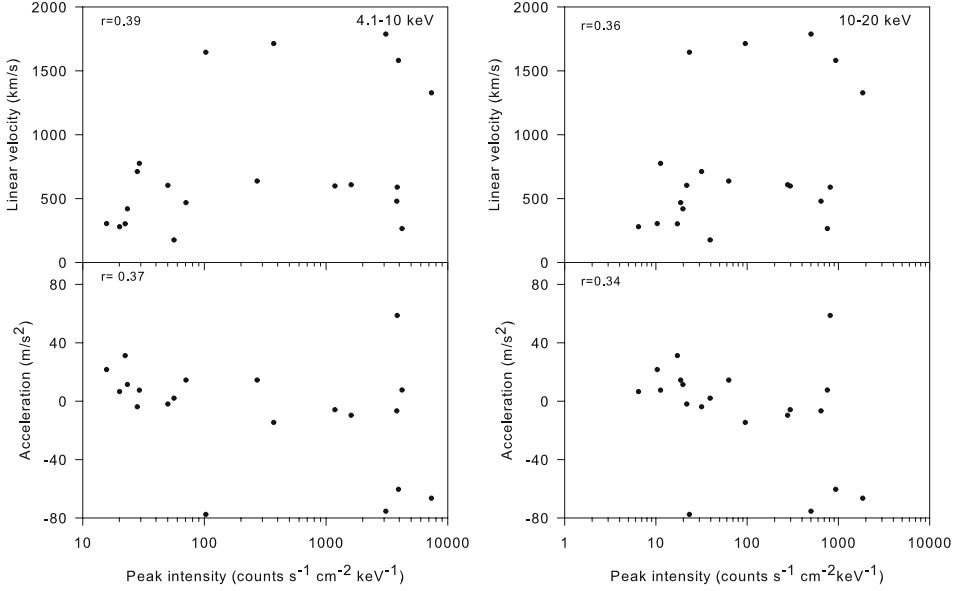


Figure 2. Velocity and acceleration of CMEs as a function of peak intensity of the flare in 4.1–10 (left) and 10–20 (right) keV energy bands.

the flux is 5σ above the background counts where σ is the root mean square of the time averaged background. End time is obtained as the time where the intensity of the flare returned to the pre-flare background or where it leveled off, whichever occurred earlier. Peak flux is the maximum flux obtained by taking the average of the measured counts in 100 ms in the flare mode and the corresponding time is the peak time of the flare. Rise time of the flare is defined as the time from the start of the flare to the peak of the emission as seen by CZT detector, while the decay time refers time from the peak of the flare to the end time. The flare duration is characterized as the time between start and end of the flare.

Shown in Fig. 2 is the variation of the linear speed and acceleration of the CMEs with peak intensity of the associated flares in 4.1–10 and 10–20 keV bands. The peak intensity of the flares varies between 15 and 8000, and 6 and 2000 counts/cm²/s/keV in 4.1–10 and 10–20 keV respectively. The CME linear speed and acceleration vary between 150 and 2000 km/s and –80 and 60 m/s² respectively. The correlation coefficient between peak intensity of the flares and linear speed and acceleration of the associated CMEs is not strong rather it ranges between 0.34 and 0.39 in both 4.1–10 and 10–20 keV energy bands.

Figure 3 represents the linear speed and acceleration of the CMEs as a function of rise time of the associated flares in 4.1–10 and 10–20 keV bands. The rise time of the flares varies between 100–900 and 50–800 seconds in 4.1–10 and 10–20 keV bands respectively. It may be noted from this figure, though not unambiguously but by and large, that speed of CMEs increases with rise time in both 4.1–10 and 10–20 keV bands, which, however, appears better in soft X-rays, viz., in 4.1–10 keV. On the other hand, we observe deceleration of CME speed with increasing rise time of the flare, which, however, appears better in 10–20 keV hard X-rays. Figure 3, therefore, indicates that fast moving CMEs showing positive acceleration are better associated with short rise

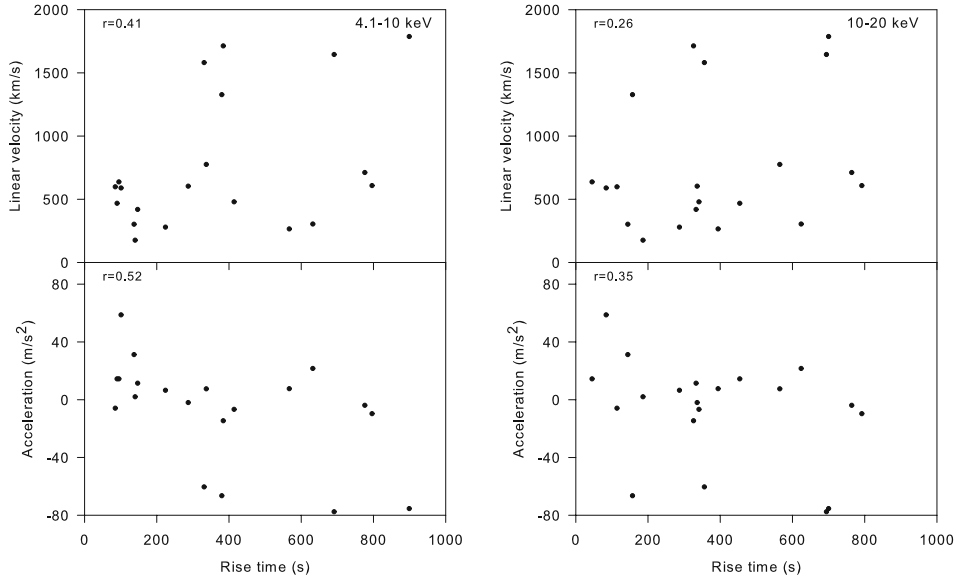


Figure 3. Variation of linear speed and acceleration of CMEs as a function of rise time of the associated flares observed with CZT detector in 4.1–10 keV (left) and in 10–20 keV (right) energy bands.

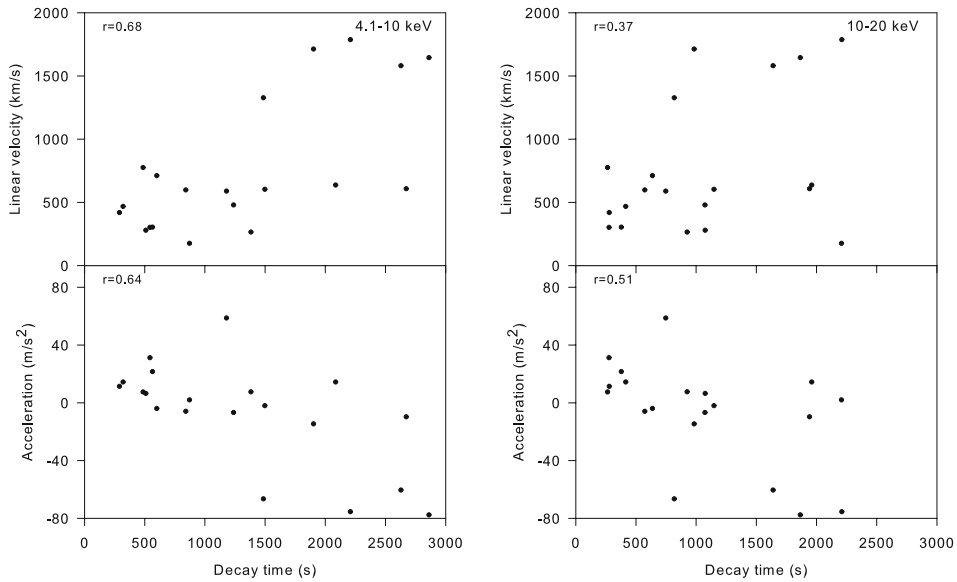


Figure 4. Variation of CME linear speed and acceleration as a function of decay time of the flare in 4.1–10 (left) and 10–20 (right) keV respectively.

time (< 150 s) i.e., impulsive flare events. However, this conclusion is subjected to limited number of events analyzed in this investigation.

Figures 4 and 5 represent the variation of linear velocity and acceleration of the CME with the decay time and duration of the associated flares respectively. We found the

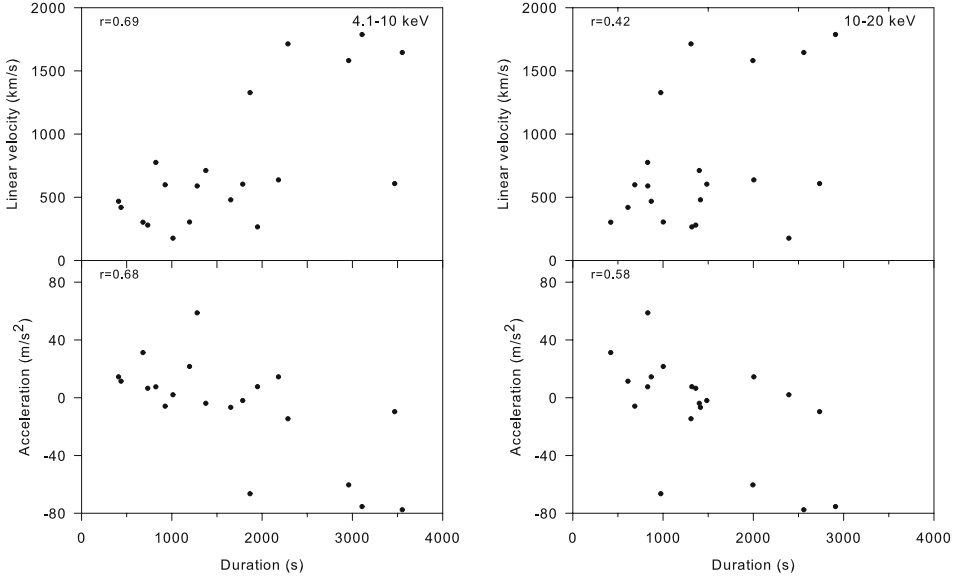


Figure 5. Variation of CME linear speed and acceleration as a function of duration of the flare in 4.1–10 (left) and 10–20 (right) keV respectively.

decay time of flare varies between 250 to 2800 and 250 to 2300 seconds in 4.1–10 and 10–20 keV energy range respectively. The higher linear velocity of CME seems to be associated with longer decay time of the flare as observed in 4.1–10 keV and 10–20 keV energy bands, however more obvious in 4.1–10 keV. This enables us to conclude that the larger speed CMEs (above 1500 km/s) are associated with flares of longer decay time (> 1500 s). However, on the contrary, CMEs decelerate faster with increasing decay time of the flares. Figure 4 also suggests that the flares associated with the CMEs of high positive acceleration decay faster than those associated with CMEs of small positive acceleration. The duration of flares (*cf.* Fig. 5) varies between 400–3550 and 400–2900 seconds in 4.1–10 and 10–20 keV respectively. We found better association of linear velocity and acceleration of CMEs with the duration of the associated flares. It may be noted from Fig. 5 that the velocity of the CMEs increases as a function of duration in both 4.1–10 and 10–20 keV bands. Again, this association appears better in 4.1–10 keV. The acceleration of CME reduces with increase of duration of flares in both the energy bands. This indicates that the possibility of CMEs with larger speeds exists with long duration flare events though they decelerate faster. However, the CMEs associated with short rise time or short duration flares show positive acceleration though they do not attain very high speeds (< 1000 km/s).

3. Discussion and conclusion

We studied temporal relationship between the dynamics of the CME and the peak intensity, rise time, decay time and duration of the associated solar flares. Our current investigation of 20 solar flare events observed by SOXS mission and found associated with CMEs suggests that higher peak intensity of the flare in 4.1–10 and 10–20 keV band is not indicative of higher speed of the associated CMEs, which is in contrast

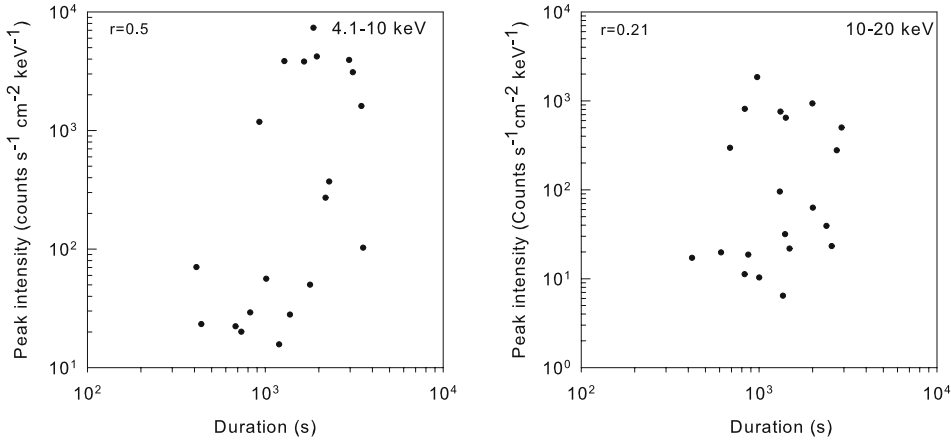


Figure 6. Variation of the peak intensity of the flares found associated with CMEs as a function of duration in 4.1–10 (left) and 10–20 (right) keV respectively.

to Munro *et al.* (1979) and Webb & Hundhausen (1987) who observed a common occurrence of the eruption and the high intensity and/or long duration flare. Kay *et al.* (2003) compared the flare parameters to find the relationship between solar flares with and without associated CMEs. They found a clear correlation between peak intensity and flare duration for the events without CMEs, while no correlation in the case of the flares found associated with CMEs, suggesting that the occurrence of a CME intrinsically affects the timescale for energy release and cooling within the flare. However, on the contrary to Kay *et al.* (2003) we found almost linear relationship between peak intensity and duration of the flare in soft X-rays, i.e., in 4.1–10 keV but not in 10–20 keV as shown in Fig. 6. This result suggests that occurrence of a CME and associated flare are two components of one energy release system that is seen at the reconnection site in the corona (Shibata 1996). Increasing soft X-ray flare intensity with duration appears to be the consequence of rise in temperature, perhaps, caused by the outgoing CME. On the other hand, no appreciable relationship between peak intensity and duration of the flare in 10–20 keV band suggests that occurrence of CME intrinsically affects the time scale for energy release in high energy bands of the flare X-ray emission, which has been described below.

We found that the faster CMEs are better associated with flares of short rise time (< 150 s) in both 4.1–10 and 10–20 keV energy bands. However, the CME associated flares showed significantly longer decay times in contrast to the rise-times, which is in agreement to Harrison (1995). The linear velocity of CMEs increases as a function of duration of flare in both 4.1–10 and 10–20 keV bands. Harrison (1995) showed that the CMEs could be associated with flares of any duration but better associated with long duration flare events, which supports our current findings. However, on the other hand, we observed deceleration in CME speed with increasing rise and decay times and duration of the flare, predominantly in hard X-ray 10–20 keV band, while CMEs associated with short rise time and/or short duration flares showed positive acceleration though they did not attain speeds higher than 589 km/s. Our current investigation shows that the linear velocity of the CME increases in general with rise time, decay time and duration of the flare, which suggests that both flare and CME are closely associated with each other and might be an integral part of the same energy release system. Further,

based on the onset time of the flares and associated CMEs within observing cadence of CMEs by LASCO/SOHO, we found that in 16 cases CME (*cf.* Table 1) preceded the flare by 23 to 1786 s, while in 4 cases the flare occurred before the CME by 47 to 685 s. Our findings support the earlier results by Harrison (1991) and more recently by Kim *et al.* (2005), and enable us to conjecture the hypothesis that flare and CME both are driven by the same magnetic activity. According to Harrison (1995) such magnetic activity may arise from shear or reconnection within the magnetic structure, and the flare and CME simply reflect the responses in different magnetic environments within the structure. More recently Zhang *et al.* (2006) proposed flux rope eruption caused by magnetic reconnection with implication in coexistent flare–CME events. They showed that an isolated flux rope coexists with two current sheets: a vertical one below and a transverse one above the flux rope. The flux rope erupts when reconnection takes place in the current sheets, and the flux rope dynamics depends on the reconnection sequence in the two current sheets. Zhang *et al.* (2006) argue that both breakout-like and tether-cutting reconnections could be important for CME eruptions and associated surface activities, which is in consistence to our findings.

We conclude that the linear velocity of CME does not depend significantly upon the peak intensity of the flare, while in general it increases with rise time, decay time and duration of the flare. Based on our findings we propose that flare and CME are the two components of one energy release system.

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Helioseismic Ring Analysis of CME Source Regions

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Abstract. We apply the ring diagram technique to source regions of halo coronal mass ejections (CMEs) to study changes in acoustic mode parameters before, during, and after the onset of CMEs. We find that CME regions associated with a low value of magnetic flux have line widths smaller than the quiet regions, implying a longer life-time for the oscillation modes. We suggest that this criterion may be used to forecast the active regions which may trigger CMEs.

Key words. Sun—coronal mass ejections—oscillations—magnetic field.

1. Introduction

Using high-resolution Global Oscillation Network Group (GONG) Dopplergrams, we have analysed the changes in oscillation mode properties in the source locations of CMEs. This was undertaken to find out if the mode properties of the CME regions are different from the active regions which did not produce a CME. This may help to identify the regions on solar disk which may produce CMEs.

2. Data and analysis

The location of the CMEs are taken from the catalogue of Zhou *et al.* (2006) who have identified the source regions of 288 halo CMEs during the period of 1997–2003. This database lists the location, date, time of occurrence and the classification of the region where the CMEs occurred. To infer the mode parameters, we use the ring-diagram analysis (Hill 1988) and compute power spectra of solar oscillations over smaller patches of the solar surface corresponding to the CME regions. The mode parameters are then extracted by fitting the power spectrum with a symmetric profile model which considers the peaks in the power spectra to be Lorentzian (Haber *et al.* 2001).

3. Results and discussion

In this analysis we examine source locations of 48 CMEs during September 2001–December 2003 that occurred no more than 30 degrees in any of the four cardinal directions from the solar disk center. These are further associated with flaring, active,

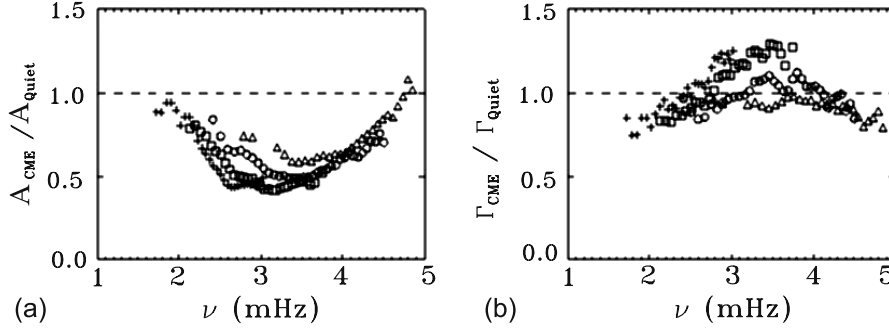


Figure 1. Changes in (a) ratio of the peak power and (b) ratio of the half-width between the CME region of December 19, 2002 ($B_{av} = 113$ G) and quiet region of December 12, 2002 ($B_{av} = 3$ G) both located at N22W12. The f -modes are represented by + and p -modes are represented by diamonds, circles, and triangles for p_1 through p_3 respectively.

and filament regions. The power spectra of these CME regions were calculated from an area covering 128×128 pixels with the center coinciding with the position of the CME. The area roughly corresponds to about 15×15 degrees in heliographic longitude and latitude. Each region was tracked for 1664 minutes where the central time corresponds to the onset time of the CME. In order to estimate the variation in mode parameters, we compare each of the CME regions with a magnetically quiet region at the same latitude as the CME region within the same Carrington rotation. The average magnetic flux (B_{av}) associated with each CME and quiet region is calculated from the synoptic maps of Kitt Peak magnetograms (<ftp://nsokp.nso.edu/kpvt/synoptic/mag/>) by averaging the unsigned magnetic flux inside the same 15×15 degree box of the corresponding regions.

As an example of the variation in mode parameters in CME regions, we have shown the ratio of the peak power (A) between a CME and a quiet region in Fig. 1(a). It is evident that f and p -modes in the CME region have lower power than the quiet region and the variation is not a monotonic function of the frequency. The maximum suppression seems to occur in the frequency range of about 3–3.5 mHz and the suppression decreases with increasing radial order. At high frequencies, the suppression tends to become small and in a few cases, we even note enhancement of power. In general, the results are consistent with those related to active regions (Rajaguru *et al.* 2001).

The variation in half-width (Γ) between the same pair of CME and quiet region is shown in Fig. 1(b) as the ratio between the two. For f and p_1 modes, the half-widths are larger in the higher end of the frequency range and smaller in the lower end of the frequency range. The widths of other p -modes are generally smaller than those in quiet regions. Since the lifetime of a mode is inversely proportional to its half-width, the small width indicates that some modes live longer in regions where CMEs originate and implies a slow damping process. This result is in contrast to the study of active regions (Rajaguru *et al.* 2001) where the lifetime of the modes were found to be smaller compared to the quiet regions.

Figure 2 shows the frequency averaged ratios of the half-width between the CME and quiet region for 48 CMEs as a function of average magnetic flux of the CME region. The widths are averaged over the frequency range of 2550–2750 μHz for f -modes and 3000–3500 μHz for p -modes. It is evident that the line width depends

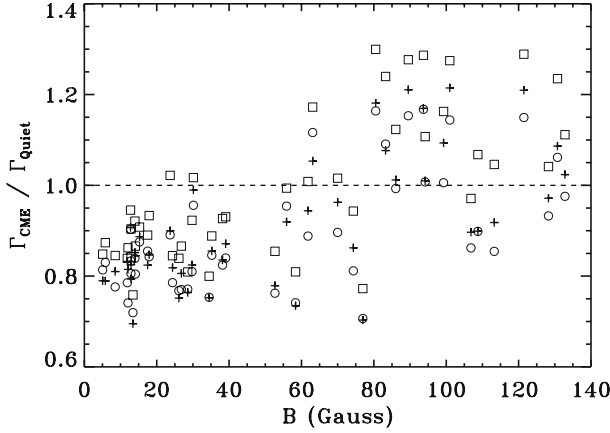


Figure 2. Frequency averaged ratio of half-widths for 48 CME regions as a function of magnetic flux. The symbols have the same meaning as in Fig. 1.

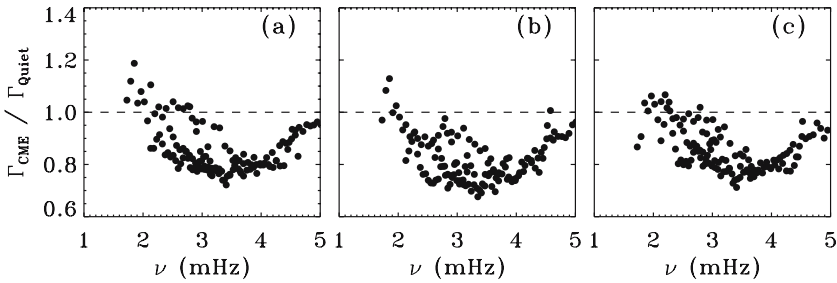


Figure 3. The changes in half-width as a ratio between a CME and quiet region for three consecutive days. Panel (b) refers to the day of onset of the CME while panels (a) and (c) refer to one day before and after the CME occurred, respectively.

on the strength of magnetic field present in the region. Most of the CMEs associated with low magnetic flux have a smaller width compared to the quiet region. The width increases for higher magnetic flux values. However, there are exceptions in both the cases.

We have also studied the evolution of the oscillation mode parameters of the CME region one day prior and one day after the onset of the CME. As an example, we show the event of March 17, 2002 where the CME is located at W10S0 and has an average flux of 14 G (Fig. 3). The quiet region for this event corresponds to the same position on April 4, 2002 and has a magnetic flux of 6.4 G. To compare, we also show the changes in half-width for three consecutive days for the active region NOAA 9628 located at S22.5E15 on September 23, 2001 (Fig. 4). On comparing Figs. 3 and 4, it is evident that in the region where the CME is triggered, the line width is always less than the quiet region, while in the case of active region, it is always higher than the quiet region. We propose this characteristic property of the oscillation modes as the prime signal to identify an active region which may produce a CME.

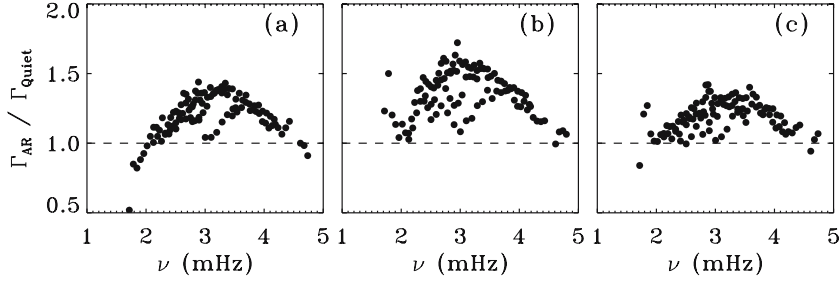


Figure 4. Same as Fig. 2 but for the active region NOAA 9628.

In summary, we find that the oscillation modes in the source locations of CMEs with a low value of magnetic flux have a higher lifetime as compared to the quiet regions. This indicates that the damping process of modes associated with CME regions are slow as compared to the quiet regions.

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Analysis of Ion Charge States in Solar Wind and CMEs

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Abstract. We discuss needs in dielectronic recombination data motivated by recent work directed at a quantitative understanding of ion charge states of various elements observed *in situ* in the solar wind and CMEs. The competing processes of ionization and recombination lead to departures from collision ionization equilibrium. The use of this as a diagnostic of acceleration and heating processes of the solar wind and CMEs is sensitive to the accuracy of the atomic rates in a way that steady state ionization equilibrium plasmas are not. The most pressing need is dielectronic recombination rates for ions Fe⁸⁺–Fe¹²⁺. These are among the dominant species observed in various regions of the solar wind and CMEs, and in remotely sensed EUV spectra.

Key words. Solar wind and CMEs—Fe ion charge states—dielectronic recombination.

1. Introduction

Interpretation of ion charge states observed *in situ* in the solar wind and also in the CMEs depends on fundamental atomic data for ionization and recombination processes. Determination of dielectronic recombination (DR) data is arguably the most complex and is plagued with most uncertainties for many-electron ions such as M-shell ions of Fe. However, there is an urgent need to determine the DR rate coefficients for these ions, especially the ions of Fe⁸⁺–Fe¹²⁺ and Fe¹⁴⁺. These are arguably the most important minor ions in solar physics, having as they do strong emission lines in the extreme ultraviolet (the bandpasses of EIS; the Extreme Ultraviolet Imaging Spectrometer, on Solar-B/Hinode were chosen to maximize the coverage of these lines), and being also the dominant charge states of Fe observed in the fast and slow solar wind. In recent years, a significant world-wide effort has revised the rate coefficients for dielectronic recombination now for all L-shell ions of all elements up to Zn, and a few beyond. For Fe, this means charge states Fe¹⁶⁺ and up. These data are now incorporated into the most recent computation of steady state ionization balance. Non-equilibrium ionization balance distributions of charge states observed in various regions of the solar wind and CMEs, are considerably less forgiving of inaccuracies in the ionization and recombination rates than are cases where ionization equilibrium can be assumed.

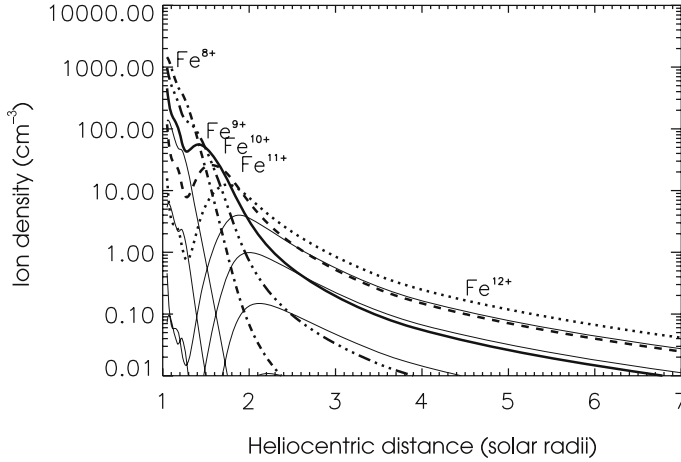


Figure 1. Plots of the evolution of the Si and Fe ion charge state densities with heliocentric distance for initial flow speed at 1.05 solar radius km s^{-1} . The initial ionization balance corresponds to the coronal hole electron temperature of 9×10^5 K. Increased ionization starts at about 1.5 solar radius, as ion-electron energy transfer increases in response to the strong ion cyclotron heating at this location. Charge states are frozen in beyond a distance of 2–2.5 solar radius, and correspond to those measured *in situ* by Ulysses.

2. Ion charge states in the solar wind: Modeling approach

The modeling of solar wind charge states generally follows the behavior of the ionization balance of a Lagrangian plasma packet, using an analytic prescription for the hydrodynamic or magnetohydrodynamic evolution. In the case of the fast solar wind, this evolution has been very well studied for polar coronal holes at solar minimum (see Laming 2004 and references therein). For halo CMEs, from which ions are detected *in situ*, we have to make use of observed or modeled height time relations for CMEs propagating across our line of sight. This situation will change with the advent of the STEREO mission. At optimum spacecraft separation, we will be able to observe directly the expansion of a CME from one spacecraft, while the CME plasma is detected at the other. Figures 1 and 2 show the evolution of Si and Fe ion charge state densities. These ion densities were obtained by solving rate equations containing electron impact ionization, radiative and dielectronic recombination rates for each charge states. These rates are the same as those used in the recent ionization balance calculations of Mazzotta *et al.* (1998), with the various updates. Dielectronic recombination from H- to He-like and from He- to Li-like are taken from Dasgupta & Whitney (2004) and DR rates for other ions were obtained from various other sources.

3. Atomic physics and ionization balance

Ionization and recombination cross sections and rates can be calculated in the Distorted Wave (DW) or Close-Coupling (CC) approximations. The DW approximation is essentially perturbation theory, and is an excellent approximation for highly charged ions, since the perturbation produced by the scattering electron is small compared to the potential acting on the target electrons due to the nuclear charge. The second approach,

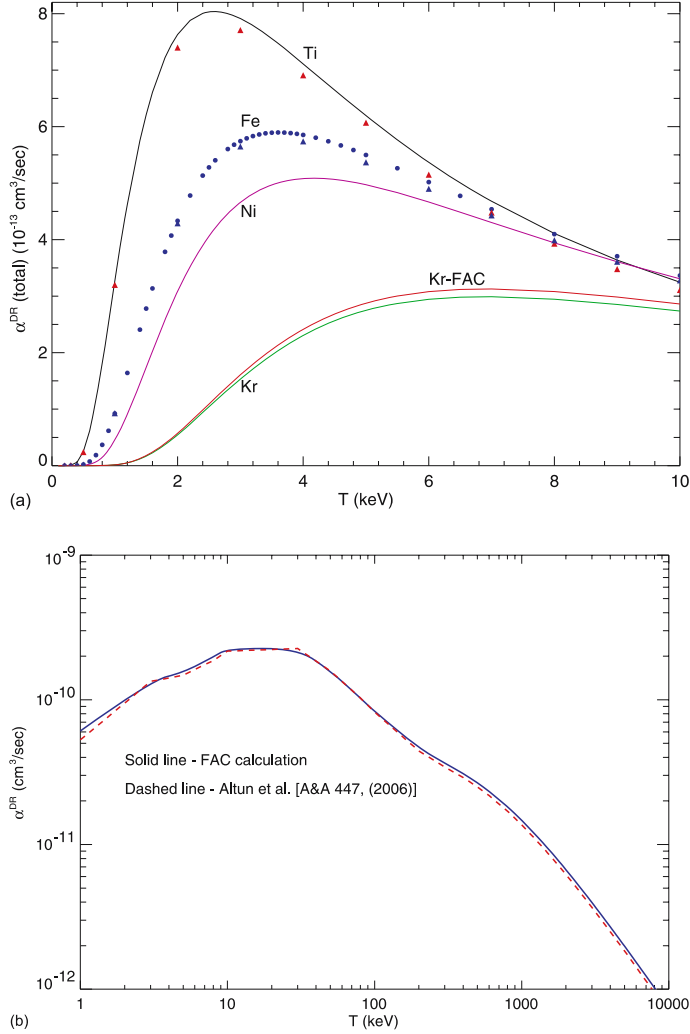


Figure 2. (a) total dielectronic recombination rate coefficients for He-like ions, Also shown are the scaled rate coefficients for Fe XXV (solid circles). The solid triangles are the results of Chen (1986a). DR rates for Kr using Cowan’s code and FAC are in good agreement. (b) dielectronic recombination rates from ground state of Fe XVI. Solid line is FAC calculations and dashed line is calculations of Altun *et al.* (2006).

CC, also known as “R-matrix” (see Berrington *et al.* 1978), entails a full integration of the Schrödinger (or Dirac) equation for the electron + target system. Resonances are included naturally so long as coupling to the resonant states is included in the wavefunction expansion. Resonances must be included “by hand” in DW approaches. Not being based on perturbation theory, R-matrix approach is suitable for less highly charged ions, but at the expense of a large increase in computation time.

We focus our investigations in calculating dielectronic recombination rates since calculations of M-shell DR rates pose major challenges due to the complexities involved

in handling the complex atomic structures of large number of excited levels for these ions. Especially relevant is the determination of charge states of ions Fe^{8+} – Fe^{12+} observed in the solar wind environment. These are important in CME cavities where Fe recombines to Fe^{8+} and beyond. In the fast solar wind Fe^{8+} – Fe^{12+} represent the dominant Fe charge states. Although there exist a series of DR data for dynamic finite-density plasmas of L-shell ions, this is not appropriate for higher temperature Fe plasmas in the solar wind and detailed and accurate calculations must be performed for reliability. However, a recent calculation by Badnell (Badnell 2006) reports high temperature DR rate coefficients for M-shell Fe ions.

Most published work based on simple formula of Burgess (1965) and other detailed calculations give only Maxwellian averaged DR rate coefficients and not DR cross sections. For studying events in the solar wind, it is extremely important to have the cross sections which can be averaged over any relevant electron distribution to generate the rates. Unfortunately there is very little to almost none of these cross sections that are available in the literature, and in their calculations of ionization balance, Laming & Lepri (2007) had to simulate rates integrated over a kappa distribution by an appropriate sum over several Maxwellian averaged rates. Maxwellian averaged rate coefficients for dielectronic recombination of course facilitate comparisons with other published work.

In Fig. 2(a), we show a comparison of DR rate coefficients of several Li-like ions obtained using Cowan's code. Detailed calculations were carried out for a few ions and then scaling laws were developed to generate data for other ions in the isoelectronic sequence. We show a comparison of DR rates obtained using Cowan's code and FAC for Kr XXXIV. FAC results agree very well with those obtained using Cowan's code.

In Fig. 2(b), we show the DR rate coefficient for Fe XVI for a wide range of temperature obtained using FAC and compare them with those obtained by Altun *et al.* (2006) using the AUTOSTRUCTURE code. In FAC, DR data are calculated in the independent-process, isolated resonance approximation. In this calculation, $\Delta N = 0$ channels that contribute mainly and significantly to the low-temperature DR rates are included along with the $\Delta N = 1$ excitations. Both $\Delta N = 0$ and $\Delta N = 1$ DR resonances of Fe XVI were measured in storage-ring experiments by Linkemann *et al.* (1995). We can compare our resonance strengths with their data by convolving with the experimental energy resolution.

Very good agreement of our previous calculations of DR rates for a number of ions with the results of complex multi-configuration fully-relativistic results of Chen for Ar through Mo, and the sample DR calculation of M-shell Fe XVI using FAC, gives us extreme confidence in obtaining accurate DR rates Fe for all the M-shell ionization stages of concern.

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The Solar Wind: Our Current Understanding and How We Got Here

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Abstract. In the original theory for the solar wind, the electron pressure gradient was the principal accelerating force. This was soon recognized to be insufficient to drive the high-speed streams. Subsequently, the discovery of Alfvén waves in the solar wind led to a long series of models in which wave pressure provided additional acceleration, but these wave-driven models ultimately failed to explain the rapid acceleration of the fast wind close to the Sun. An alternate view was that the pressure of hot protons close to the Sun could explain the rapid acceleration, with the proton heating coming from the cyclotron resonance. SOHO has provided remarkable data which have verified some of the predictions of this view, and given impetus to ongoing studies of the ion-cyclotron resonance in the fast wind. After a historical review, we discuss the basic ideas behind current research, emphasizing the importance of particle kinetics. We conclude with some guesses as to how work might proceed in the future.

Key words. Solar wind—coronal holes—solar corona—cyclotron resonance—MHD waves.

1. Introduction

Early spacecraft data in the 1960s revealed solar wind properties which could not be well-explained by models in which the electron pressure gradient was the principal accelerating force. The Alfvén waves discovered around 1970 to be plentiful in the solar wind were thought for a while to provide additional energy and momentum, but they ultimately failed to explain the rapid acceleration of the fast wind close to the Sun. The issue remained clouded until the mid-1990s when the Solar and Heliospheric Observatory (SOHO) began yielding remarkable data concerning the origin of the solar wind, and the heating of the corona from which it originates. (The coronal heating problem is certainly the most famous, but it is often forgotten that the power requirements of the chromosphere per unit of solar surface area are comparable to, and sometimes greater than those of the corona.) SOHO has forced us to discard some long-held ideas, and has also stimulated a number of entirely new ideas. The purpose of this review is to survey our current understanding of the fast solar wind, emphasizing the particle kinetics and the roles of the cyclotron resonance. We also emphasize that there are many remaining problems, especially the source of the ion-cyclotron waves. Since we may be at a watershed, we will also review the historical context.

We will, for the most part, concentrate on the fast solar wind for four reasons: First, we have a pretty good idea where it comes from, especially the large polar coronal holes; see Miralles *et al.* (2002, 2004) for discussions of solar wind flows out of smaller non-polar holes. Second, the fast wind is much steadier than the slow wind, and seems to be less structured. It has always been hoped that this apparently simpler wind would be more amenable to yielding its secrets. But in spite of this, the waves and turbulence in the slow wind seem to have less power than in the fast wind (e.g., Tu & Marsch 1995). Third and most importantly, the fast wind has long been known to be less influenced by Coulomb collisions than the slow wind (e.g., Neugebauer 1981). This means that signatures of other kinetic processes are more readily seen in the data. Indeed, one of our conclusions will be that microscale kinetic processes are an essential ingredient of the physics of the fast wind. Fourth, the physics of the slow wind is widely believed to be very different from that of the fast wind. Thus the slow wind really requires an extensive review of its own; see for example Cranmer (2005); Einaudi *et al.* (1999); Wang *et al.* (2000); Fisk & Schwadron (2001); Lapenta & Knoll (2005); and references therein.

The reader should be aware that we are not attempting a complete literature survey. We do try, however, to refer key papers which themselves provide references to the bulk of the literature. See especially recent reviews by Hollweg & Isenberg (2002, hereafter ‘HI’); Cranmer (2002, 2004); Erdös (2003); and Hollweg (2006a, 2006b). We also recommend the review of the physics of wave-particle interactions by Tsurutani & Lakhina (1997).

2. The electron-driven wind

Parker (1958) noted the thermal conductivity of ionized hydrogen scales as $T_e^{5/2}$, where T is temperature and subscript e denotes electrons. If outward thermal conduction dominates the heat equation, then $T_e \propto r^{-2/7}$ in a spherically-symmetric corona, r being heliocentric distance. For a corona in static equilibrium, Parker showed that the slow decline of T_e leads to a plasma pressure at $r \rightarrow \infty$ which is many orders-of-magnitude larger than the interstellar pressure. He concluded that with nothing to contain the asymptotic pressure, the corona must expand; the expansion is driven by the electron pressure gradient because T_e remains high.

However, in a later review, Parker (1965) compared theoretical predictions with the by then known properties of the solar wind at 1 AU. He concluded “thus the model for the hypothetical conduction corona leads to a temperature falling too rapidly with radial distance from the Sun, indicating that the actual solar corona is probably actively heated for some considerable distance by the dissipation of waves”. We are not yet sure whether waves are responsible, but we now know that he was right about extended coronal heating.

Hartle & Sturrock (1968) presented the first two-fluid model of the solar wind, with separate energy equations for electrons and protons. They obtained a very slow wind: 250 km s^{-1} at 1 AU compared to a typical fast wind speed of 750 km s^{-1} (see Feldman *et al.* (1976) for an early review of the properties of the fast wind). They also found that the model protons were much too cold: 4400 K at 1 AU compared to several times 10^5 K in fast wind. They concluded that “departures of the solar wind characteristics near Earth from those of this model are also to be attributed to heating by a flux of non-thermal energy”.

Neither Parker nor Hartle and Sturrock said anything about coronal heating. They began their models at $r = r_S$ with 10^6 K coronal temperatures, and no subsequent heat addition (r_S is the solar radius).

3. The wave-driven wind

The next major advance was the discovery (Belcher & Davis 1971) of the ubiquitous presence of Alfvén waves in the solar wind. Most of the wave power resides at long periods, of the order of hours. The waves predominantly propagate away from the Sun, especially in the fast wind. The outward propagation strongly suggests that the Sun is the source of these waves. One of the outstanding questions is: what on the Sun is responsible for the dominance of the observed timescales of hours?

It was immediately realized that the Alfvén waves might be Parker’s “waves”, and Hartle and Sturrock’s “flux of non-thermal energy”. Alazraki & Couturier (1971) and Belcher (1971) inaugurated the concept of the wave-driven wind by noting that the waves exert a ‘wave pressure’ $-\nabla \langle \delta \mathbf{B}^2 \rangle / 8\pi$ on the wind (in cgs units, which we shall use throughout); $\mathbf{B}$ is magnetic field, the prefix δ denotes a fluctuation, and the angle brackets denote a time-average. Hollweg (1973) showed how the wave energy equation could be extended to include dissipation and plasma heating. With heating and wave pressure, the wave-driven models were able to explain the high-speeds and hot protons observed in the fast wind in interplanetary space (e.g., Hollweg 1978).

An early mention of the radiation pressure of Alfvén waves can be found in Bretherton & Garrett (1969). Their work was expanded upon by Dewar (1970) and Jacques (1977).

These wave-driven models generally succeeded in explaining solar wind data far from the Sun, but they failed close to the Sun. In the early 1990s, spacecraft gave us new coronal hole density data, which verified previous evidence that the density declines very rapidly with increasing r (Guhathakurta & Holzer 1994; Fisher & Guhathakurta 1995; Guhathakurta & Fisher 1998). That requires the flow speed to increase very rapidly with r . The wave-driven models could not achieve such rapid accelerations. The reason is simply that, close to the Sun, the wave pressure is small compared to other terms in the momentum balance.

There were other difficulties as well, particularly concerning heavy ions, the best observed being He^{++} because of its relatively high abundance. Ions flow faster than the protons, roughly by the Alfvén speed V_A , and they are hotter than the protons, roughly in proportion to their masses; see Neugebauer (1992) for a review. These properties are most noticeable in the fast wind, because Coulomb collisions are weaker there. Efforts (mainly in the early 1980s) to explain these observations generally invoked the ion-cyclotron resonance to heat and accelerate the ions; see reviews by Isenberg (1983); Cranmer (2002, 2004), and HI. These models were only partially successful. One difficulty was that these early studies considered resonant effects only far from the Sun, well beyond the acceleration region. In contrast, the SOHO data show extensive heavy ion heating in the acceleration region.

There were other indications that the ion-cyclotron resonance was at work. As reviewed by Marsch (1991), *in situ* measurements of proton distribution functions often show that the protons in the vicinity of the peak of the distribution are anisotropic, with more thermal energy perpendicular to the local magnetic field than along the field; this is most noticeable in high-speed wind. Moreover, the average magnetic

moment, $T_{p\perp}/B$, increases with distance from the Sun ($T_{p\perp}$ is the proton temperature perpendicular to $\mathbf{B}$). Both of these observations suggest that perpendicular heating is occurring in interplanetary space, and that in turn suggests the cyclotron resonance.

4. The proton-driven wind

Hollweg (1986) and Hollweg & Johnson (1988, hereafter ‘HJ’) applied these ideas to the solar wind close to the Sun. They assumed that the Sun launches low-frequency Alfvén waves which undergo a turbulent cascade to high frequencies, where they are dissipated into heat via the cyclotron resonance; HJ assumed that the resonant dissipation would heat only the protons, and that this was the only source of coronal heating. The heating rate was dictated by the rate at which energy cascades to high frequencies; HJ took the Kolmogorov rate, $Q = \rho \langle \delta \mathbf{V}^2 \rangle^{3/2} / L_{\text{corr}, \perp}$ (Q is the volumetric heating rate, ρ is plasma density, $\langle \delta \mathbf{V}^2 \rangle$ is the velocity variance associated with the waves, and $L_{\text{corr}, \perp}$ is the correlation length perpendicular to the magnetic field). These models also included acceleration by the wave pressure. The waves were taken to be outward-propagating in the short-wavelength WKB limit. (We shall later see that this is formally inconsistent with turbulence.) These models succeeded in reproducing the observed high-speed wind far from the Sun, as well as the rapid acceleration close to the Sun. HJ found that the protons close to the Sun in $r > 3r_S$ were considerably hotter than the electrons. The pressure of the hot protons was mainly responsible for the rapid flow acceleration. The available data at the time indicated that the protons were not hot close to the Sun, and so these models were discarded. Isenberg (1990) extended HJ by including He^{++} . This model too had hot protons close to the Sun, and was discarded.

As it happened, HJ and Isenberg (1990) actually predicted the results which SOHO would soon obtain for coronal holes: hot coronal protons, and heavy ions which flow faster and are more than mass-proportionally hotter than the protons close to the Sun. Moreover, the SOHO results, especially from the Ultraviolet Coronagraph Spectrometer (UVCS), indicate that positive particles in coronal holes close to the Sun are heated mainly perpendicular to the magnetic field; this result is firm for O^{+5} (Doderer *et al.* 1998; Kohl *et al.* 1998; Antonucci *et al.* 2000), but less certain for protons. Thus the SOHO/UVCS results suggest:

1. Consistent with the perpendicular heating of O^{+5} , the cyclotron resonance is at work.
2. With protons hotter than electrons, the solar wind flow is mainly proton-driven, via the divergence of the (presumably anisotropic) proton pressure tensor.
3. At least in the fast wind, coronal heating is proton and ion-dominated. It is not Joule dissipation.
4. The peculiar properties of heavy ions originate close to the Sun in the wind’s acceleration region.
5. Coronal heating extends through the acceleration region into the supersonic wind. Most remarkable is O^{+5} , which has a temperature 3×10^8 K at $r = 3.5r_S$; the temperature is still an increasing function of r , in spite of strong adiabatic cooling (Kohl *et al.* 1998). Coronal heating and solar wind acceleration must be treated together. Parker (1965) was right about extended heating.
6. Other instruments on SOHO, especially SUMER (Solar Ultraviolet Measurements of Emitted Radiation), also show indications of ion heating low in the corona (e.g., Tu *et al.* 1998; Moran 2003; Peter & Vocks 2003).

There is, however, one note of caution. The proton and oxygen temperatures are derived from spectral linewidths, assuming that non-thermal broadening due to waves or turbulence is negligible. This is almost certainly true for O^{+5} , which shows considerably higher linewidths than the protons (Esser *et al.* 1999; Cranmer 2004). (Recall that $\mathbf{E} \times \mathbf{B}$ drifts would be the same for all particles if their flow speeds differ by much less than V_A ; $\mathbf{E}$ is the electric field.) Moreover, models (e.g., HI) suggest that waves or turbulence do not contribute substantially to any of the observed linewidths.

Finally, we note that other authors investigated the effects of hot coronal protons (Esser & Habbal 1995, 1996; Hansteen & Leer 1995; Axford & McKenzie 1996; Esser *et al.* 1997; McKenzie *et al.* 1995, 1997; Lie-Svendsen *et al.* 2001), but only *ad hoc* heating functions were used. Li (1999, 2002); Li & Habbal (1999); and Li *et al.* (1999, 2004) have extended the original models of HJ and Isenberg (1990) to include thermal anisotropy, wave dispersion, and other effects.

5. The ion-cyclotron resonance

The cyclotron resonance occurs when the wave frequency seen by a particle matches the particle's cyclotron frequency. Formally the resonance condition is $\omega - k_{\parallel} V_{\parallel} = \pm \Omega$, where ω is the wave angular frequency, $k_{\parallel}$ is the wavenumber along $\mathbf{B}$, $V_{\parallel}$ is the particle's drift speed along $\mathbf{B}$, and Ω is the gyro-frequency. For ions resonating with left-hand polarized waves the $+$ sign is appropriate, while the $-$ sign is used for right-hand waves. When the resonance condition is satisfied, the particle's energy changes secularly. The secular energy change can be a gain or a loss, depending on the phase of the particle's gyro-motion relative to the wave. In a random field, that phase will be random, the particle will gain or lose energy randomly, and consequently the particle will undergo a random walk in velocity space, which in turn leads to velocity space diffusion, frequently referred to as pitch-angle diffusion; heating and acceleration of a particle distribution is formally analyzed using diffusion equations in velocity space.

Not all particles can resonate. The heavy curve in Fig. 1 shows the dispersion relation for the electromagnetic ion-cyclotron wave propagating along $\mathbf{B}$ in a cold electron-proton plasma: $k^2 V_{Ap}^2 = \Omega_p \omega^2 / (\Omega_p - \omega)$, where V_{Ap} is the Alfvén speed based on the proton density (we have taken $m_e = 0$ for these low-frequency waves), and Ω_p is proton gyro-frequency; Fig. 1 is in the frame moving with the bulk proton flow. In Fig. 1, the resonance condition is a straight line with slope $V_{\parallel} / V_{Ap}$ and an ordinate intercept at Ω / Ω_p . Near the top of the figure we show the resonance conditions for two test protons, one (line 'A') with $V_{\parallel} > 0$ and moving with the wave, and the other (line 'B') with $V_{\parallel} < 0$ and moving against the wave. Resonance can only occur if the resonance condition intersects the dispersion relation. For line 'A' there is no resonance. Thus only protons moving against the wave can resonate. Also shown in Fig. 1 are resonance conditions for four O^{+5} test particles, two with $V_{\parallel} > 0$ (lines 'C' and 'D') and two with $V_{\parallel} < 0$ (lines 'E' and 'F'). Lines 'D'–'F' show that resonance is possible for oxygen ions moving both with and against the wave. In resonant heating and acceleration, this gives the ions a major advantage over the protons. However, line 'C' also shows that O^{+5} can drop out of resonance if $V_{\parallel}$ becomes too large; we will return to this point below. (Actually, He^{++} is sufficiently abundant to modify the dispersion relation. As discussed by Hollweg (2000a) and HI, He^{++} suffers a disadvantage similar to that of the protons. UVCS gives no information about He^{++} and we will ignore this issue here.)

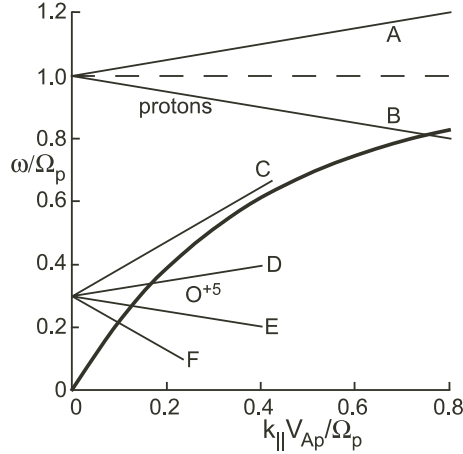


Figure 1. Dispersion relation (heavy curve) for the parallel-propagating ion-cyclotron mode in a cold electron-proton plasma, as viewed in the bulk proton frame. The resonance condition is shown as the thin straight lines, with positive (negative) slopes for test particles moving with (against) the wave. The top two lines are the resonance conditions for protons; only protons moving against the wave can be in resonance. The lower four lines are for O^{+5} ; ions moving with and against the wave can be in resonance, but they can drop out of resonance if they move too rapidly with the wave. Lines ‘E’ and ‘F’ show that O^{+5} ions moving against the wave never drop out of resonance. In fact, as they move more rapidly against the wave, they become more resonant in the sense of resonating with lower frequency waves which presumably have more power. In the corona, antisunward moving ions may be affected mainly by sunward-propagating waves.

Figure 2 illustrates several effects associated with velocity space diffusion. To keep things simple, we will here ignore dispersion, so that we can define a wave frame. Since there is no wave electric field in that frame, particles will conserve energy and diffuse along circular arcs $V_{||}^2 + V_{\perp}^2 = \text{constant}$ ($V_{\perp}$ is the velocity perpendicular to **B**). In Fig. 2, the circular arcs are centered on the phase speed in the proton frame. In reality protons will resonate with slower-moving waves than O^{+5} , so we have drawn separate circles for those particles. Marsch & Tu (2001) have actually found evidence that solar wind protons diffuse along circular arcs; they write “direct observational evidence from Helios plasma data is shown for the occurrence of this pitch angle diffusion of solar wind protons, induced by resonance with parallel ion-cyclotron waves propagating away from the Sun . . . parts of the isodensity contours in velocity space are well outlined by a sequence of segments of circles centered at the adapted wave phase speed . . . ”

We consider what happens to a group of particles that start out on the abscissa with $V_{\perp} = 0$. Those particles will diffuse upward along their arcs, thereby acquiring a perpendicular temperature; this is the most important part of the resonant heating since particles with large $V_{\perp}$ can subsequently be accelerated via the magnetic mirror force. Particles which start with the same $V_{||}$ will acquire a spread in $V_{||}$, and thus an increased parallel temperature, as they diffuse. (It is also possible for $T_{||}$ to decrease (Dusenbery & Hollweg 1981; Li *et al.* 1999). $T_{||}$ heating or cooling is not a major effect, and we will not consider it further.) As particles diffuse upward along their arcs, they also move to the right. This represents a bulk resonant acceleration of the particles, but it is generally not as important as the magnetic mirroring. Now note the thick portions

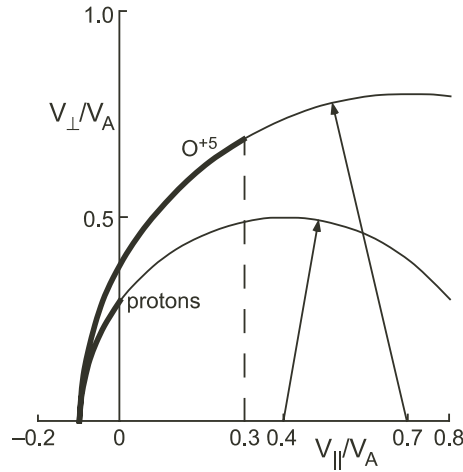


Figure 2. Particle diffusion, as viewed in the bulk proton frame. In the wave frame particle energy is conserved, so particles diffuse along circular arcs centered on the wave phase speed. The arcs for O^{+5} and protons are centered on different values of $V_{\parallel}$, because O^{+5} tends to resonate with faster-moving waves than do the protons. The thicker portions of the arcs indicate the values of $V_{\parallel}$ for which the particles can be in resonance. O^{+5} (and other heavier ions as well) can diffuse to higher values of $V_{\perp}$ than can the protons. Thus if the diffusion is fast enough, ions can become more than mass-proportionally hotter than the protons.

of the arcs in Fig. 2. They indicate the values of $V_{\parallel}$ for which particles can be in resonance. Only protons with $V_{\parallel} < 0$ (i.e., moving against the wave) are in resonance, while O^{+5} ions drop out of resonance only when $V_{\parallel}$ is sufficiently positive (i.e., moving with the wave). Diffusion will tend to fill up the dark arcs, but not beyond. This fact, along with the two circles having different radii, clearly allow O^{+5} to attain larger values of $V_{\perp}$ than the protons. Recalling that temperature is proportional to mass times (thermal speed)², we see that diffusion tends to give $T_{\text{oxygen}, \perp} / T_{p \perp} > m_{\text{oxygen}} / m_p$, in qualitative agreement with the UVCS/SOHO results (m denotes particle mass). Finally, the oxygen advantage is furthered by the fact that O^{+5} resonates with lower frequency waves which are observed to have more power.

These considerations are qualitatively in accord with the UVCS/SOHO data, but there are quantitative difficulties. Kohl *et al.* (1998) show that the O^{+5} temperature is an increasing function of r in $2 < r/r_S < 3.5$ (the outer limit of the data). But models (Hollweg 2000a; HI) involving resonances with outward-propagating waves have been unable to reproduce this result. The resonant heating is effective in producing very hot O^{+5} rather close to the Sun. But the large values of $V_{\perp}$ then lead to a rapid acceleration via the mirror force, causing the oxygen to drop out of resonance, and to experience strong adiabatic cooling.

Resonances with sunward-propagating waves can save the day. (Sunward-propagating waves can arise from reflections or instabilities.) An outward-moving ion resonating with a sunward-propagating wave will have a resonance condition with negative slope in Fig. 1, e.g., line ‘E’ in the figure. As the particle accelerates outward while still interacting with inward waves, the slope of the straight line will become even more negative, e.g., line ‘F’. Not only will the particle never drop out of resonance, but it will resonate at ever smaller wavenumbers, where there is more power.

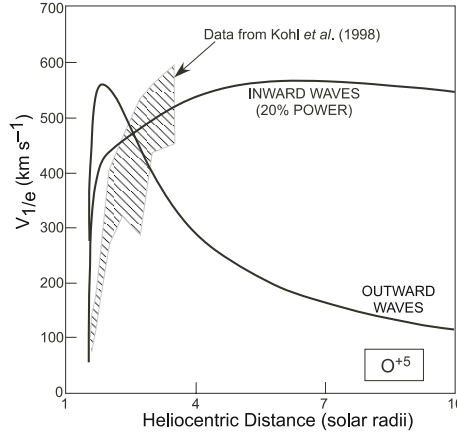


Figure 3. The thermal speed $V_{1/e}$ of O^{+5} in a simple model which demonstrates that the effects of inward- and outward-propagating waves are qualitatively different, for the reasons discussed in the text. The cross-hatched area represents the coronal hole data from Kohl *et al.* (1998). These data are remarkable in that they show that the O^{+5} temperature increases with increasing r out to $3.5r_s$ (the outer limit of the data), in spite of strong adiabatic cooling. Outward-propagating waves fail to explain this, even qualitatively, because the particles drop out of resonance. Inward-propagating waves do not suffer this defect, and generally succeed in matching the data.

The efficacy of this idea has been illustrated in a simple “proof of principle” model of the acceleration of O^{+5} in a coronal hole (Hollweg 2006c). In Fig. 3, we show the perpendicular heating of O^{+5} (as represented by the perpendicular thermal velocity) for two cases: pure outward-propagating waves, and pure inward-propagating waves with one-fifth the power of the outgoing waves. The two cases produce very different behaviors, even qualitatively. Particles resonating with outgoing waves are heated very rapidly close to the Sun, thereby acquiring large $V_{\perp}$ which results in a strong outward mirror acceleration, which in turn causes the particles to drop out of resonance (line ‘C’ in Fig. 2) and cool adiabatically at greater distances from the Sun. In contrast, particles resonating with ingoing waves never drop out of resonance, and have $V_{\perp}$ increasing with r in regions where particles resonating with outgoing waves have $V_{\perp}$ decreasing with increasing r . The grey shading in Fig. 3 indicates the UVCS data from Kohl *et al.* (1998); values of $V_{\perp}$ for particles resonating with outgoing waves do not resemble the data at all, while particles resonating with inward waves show a better match. Hollweg (2006c) also finds that the observed behavior of the bulk flow speed of O^{+5} is better explained by interactions with ingoing waves. Detailed modeling of the ion resonances will have to carefully consider the mix of sunward-propagating and antisunward-propagating waves. Unfortunately, this represents a significant lacuna in our understanding, since we are not sure where the ion-cyclotron waves come from in the first place.

Another effect associated with sunward-propagating waves has been discussed by Isenberg (2001a), who pointed out that O^{+5} can simultaneously resonate with sunward and antisunward waves, while the protons cannot. (To see this, imagine the dark arcs in Fig. 2 flipped about the vertical axis. Those flipped arcs will represent pitch-angle diffusion caused by waves propagating to the left in the figure. Now imagine the

flipped arcs superimposed on the arcs that are already in the figure. The two sets of arcs will represent the pitch-angle scattering which occurs when waves propagating both to the left and to the right are present. The two sets of arcs for protons will not overlap; this means that a proton can resonate either with a left-propagating wave or with a right-propagating wave, but not with both waves simultaneously. In contrast, the two sets of arcs for O^{+5} will overlap to some extent; this means that a given O^{+5} ion can resonate with left- and right-propagating waves simultaneously.) In terms of diffusion, protons diffuse only along the arcs, while O^{+5} can diffuse across the arcs. This cross-arc transport is more commonly called second-order Fermi acceleration (e.g., Terasawa 1989). Since it is available to O^{+5} but not to the protons, it represents an energization mechanism which is inherently preferential to heavy ions. (He^{++} would tend to behave like protons in this scenario.) Quantitative evaluation of the simultaneous effects of sunward- and antisunward-propagating waves remains to be carried through. (See Isenberg & Vasquez (2006) for some early fully kinetic computations of how distribution functions evolve when resonances with both ingoing and outgoing waves are operative, but it should be kept in mind that, that paper does not include the tendency of the mirror force to accelerate the ions to the point where they drop out of resonance with outgoing waves. The author strongly suspects that once the ions are accelerated outward by the mirror force, they will interact mainly with the sunward-propagating waves.)

Another puzzle concerns Mg^{+9} , which has also been studied by SOHO/UVCS. Whereas O^{+5} is heated more than mass-proportionally relative to the protons, Mg^{+9} attains temperatures which are roughly mass-proportional. Why is there such different behavior for ions with similar values of q/m ? The answer is not known; see discussions in HI and Cranmer (2002).

6. Whence the ion-cyclotron waves?

The UVCS/SOHO data seem to be plangent evidence for the cyclotron resonance. But there is no agreement on where the ion-cyclotron waves come from.

Based on seminal work by Coleman (1968) and Barnes (1979), one school of thought follows Hollweg (1986) and HJ: the Sun launches low-frequency waves which undergo a turbulent cascade that produces the high-frequency resonant waves. This viewpoint is well-motivated observationally. *In situ* observations of magnetic field and velocity fluctuations show most power at low frequencies, but with power law power spectra extending to high frequencies (e.g., Marsch & Tu 1990). The power law indices have a preference for the $-5/3$ value expected for a Kolmogorov turbulent cascade. *In situ* data for the corona are not available. Radio studies do give some information about density fluctuations, though their connection with ion-cyclotron waves is unclear. The data are suggestive nonetheless. Most power resides at low frequencies (corresponding to periods of hours), and the power spectra have the Kolmogorov index at higher frequencies, with some flattening just below the turbulence dissipation range (Coles & Harmon 1989). In both radio and *in situ* data, the power spectra steepen at the spatial scales expected if the ion-cyclotron resonance is coming into play (e.g., Leamon *et al.* 1998a, 1998b, 1999, 2000; Yamauchi *et al.* 1998; Bale *et al.* 2005; Harmon & Coles 2005); the spectral steepening is identified with the dissipation range.

It is often suggested that the turbulence scenario has a serious flaw: MHD turbulence tends to produce large cross-field wavenumbers, $k_{\perp}$, rather than the large values of

$k_{\parallel}$ needed for cyclotron resonance (e.g., Shebalin *et al.* 1983; Ng & Bhattacharjee 1996; Milano *et al.* 2001; Oughton *et al.* 2004). But the observational fact is that there is substantial power at large $k_{\parallel}$'s. This is especially true in the high-speed wind at low frequencies (Dasso *et al.* 2005), but also near the dissipation range (Leamon *et al.* 1998a, 1998b, 1999, 2000). From radio studies close to the Sun, Harmon & Coles (2005) have concluded "radio scattering indicates there must be a substantial parallel component to the wave power which, in the turbulence picture, suggests a substantial non-perpendicular cascade". Vasquez *et al.* (2004) have proposed that large $k_{\parallel}$'s can be produced by turbulence if the background is spatially structured; the waves advect the structures, which in turn refract the waves, and so on, quickly leading to large $k_{\parallel}$'s. Perhaps non-MHD processes are at work, particularly near the dissipation range where kinetic effects can come into play. We believe that a particularly important point has been made by Chandran (2005) who writes "incompressible and weakly compressible MHD neglect the fast wave . . . when fast waves are accounted for, MHD turbulence in low- β plasmas transfers energy to high-frequency fast waves and, to a lesser extent, high-frequency Alfvén waves". In other words, the common wisdom that MHD turbulence only produces high $k_{\perp}$ may be wrong. (See also Cranmer & van Ballegoijen (2003) for a discussion of the $k_{\parallel}$ problem.)

If the turbulence scenario is correct, the Sun must launch sufficient power at long periods to drive the fast wind; an energy flux density of $5 \times 10^5 \text{ erg cm}^{-2} \text{ s}^{-1}$ at the coronal base is a representative number. This issue has been addressed by looking at Faraday rotation fluctuations impressed on a radio signal as it traverses the corona. Hollweg *et al.* (1982) used the linearly polarized signal from the Helios spacecraft. They found that the polarization direction varied with timescales of the order of hours, suggesting a connection with the long-period Alfvén waves observed *in situ*. They showed that the observed Faraday rotation fluctuations in $2r_S < r < 15r_S$ closely matched what would be expected if there were indeed long-period Alfvén waves in the corona with enough power to drive the fast solar wind. Similar results were obtained by Andreev *et al.* (1997). However, Mancuso & Spangler (1999) and Spangler (2002) looked at Faraday rotation fluctuations from natural radio sources and concluded that there was not enough long-period wave power to drive the fast wind. But it should be pointed out that there are two basic problems with the Faraday rotation studies. The first is that the observed rms fluctuations depend on the line-of-sight correlation length for the turbulence, which has to be guessed. The second difficulty is that the studies so far have extrapolated the data back to the coronal base without allowing for dissipation; if these waves ultimately heat the corona and drive the fast wind, there has to be significant dissipation. (For example, HI found that the wave action falls off approximately as r^{-1} in a simple model with turbulent dissipation.) Allowing for dissipation would lead to a larger value of the energy flux density at the coronal base.

Harmon & Coles (2005) have used other radio data, sensitive to the density fluctuations, to investigate the energetics of the turbulence cascade close to the Sun. Their conclusion is "the cascade energy dissipated in proton-cyclotron damping and electron Landau damping is large enough to be an important factor in solar wind heating and dynamics".

Like the *in situ* data, the Faraday rotation and density fluctuations in the corona usually have timescales of the order of hours (not the 5-minute or 3-minute periods we are most accustomed to hearing about, though we will mention an exception in the following paragraph). What on the Sun is responsible for these timescales? We know of

only one solar phenomenon which occurs on timescales of hours: the flux cancellation events (Livi *et al.* 1985; Martin *et al.* 1985; Ryutova *et al.* 2003). It is not unreasonable to suppose that the drastic alteration of magnetic field in a flux cancellation event could launch long-period Alfvén waves with substantial energy fluxes (Hollweg 1990). This conclusion is probably closely related to recent studies by Close *et al.* (2004, 2005). They found that “the timescale for magnetic flux to be remapped in the quiet-Sun corona is, surprisingly, only 1.4 hr . . . implying that the quiet-Sun corona is far more dynamic than previously thought.” In view of the long-known presence of hour periods in the low corona and solar wind, they shouldn’t have been surprised. (Twisting or shaking of intense magnetic flux tubes by the convective motions should also launch Alfvén waves. The trouble is that the timescales aren’t right. The solar granulation has timescales of minutes while the supergranulation has timescales of tens of hours. The photosphere also contains weaker fields, called inter-network or intra-network fields (e.g., Lites & Socas-Navarro 2004; Cerdeña *et al.* 2006); the possible role of these fields in launching waves has, to our knowledge, not been explored.)

A variation of the turbulence scenario is suggested by results of Ulrich (1996). He looked at line-of-sight velocity and magnetic field fluctuations in the chromosphere, and found that they were correlated consistent with outgoing Alfvén waves. The time-averaged upward Poynting flux was about the amount required to power the chromosphere and corona in strong field active regions, where the data were taken. The study has not been done, but it is not unreasonable to suppose that the Alfvénic energy flux in regions of weaker fields, e.g., in coronal holes, might be sufficient to drive the high-speed solar wind. However, unlike the *in situ* data and much of the radio data, Ulrich’s fluctuations were mainly in the 5-minute band. In this context, it is interesting to note that Chashei *et al.* (1999) have provided observations of Faraday rotation fluctuations (due to magnetic field fluctuations which are believed to be associated with Alfvén waves) of radio signals propagating through the corona in the acceleration region of the solar wind; there is evidence of a peak at periods around 5 minutes, in an otherwise broad power spectrum. Might the putative turbulent cascade originate from waves with periods of the order of 5 minutes, rather than hours? (See also Cranmer & van Ballegooijen (2005) for a discussion of Alfvén wave generation by random walks, induced by the solar “5-minute oscillations”, of photospheric magnetic flux tubes.)

An alternate proposal for the origin of the ion-cyclotron waves is that magnetic reconnection events in the photosphere or chromosphere launch waves with the kilohertz frequencies that are resonant with protons and ions in the corona (e.g., Schwartz *et al.* 1981; Axford & McKenzie 1992, 1996; McKenzie *et al.* 1995, 1997; Czechowski *et al.* 1998; Ruzmaikin & Berger 1998). To generate kilohertz frequencies, the reconnecting elements would have to be extremely small. The waves originate well below the local ion-cyclotron frequencies in the photosphere and chromosphere, but they become resonant in the corona where the magnetic field is much weaker. Tu & Marsch (1997) and Marsch & Tu (1997) presented detailed solar wind models based on this scenario. They assumed that the waves propagate along $\mathbf{B}$ with a prescribed power spectrum. Hollweg (2000b) suggested that oblique propagation is far more likely, in which case the waves would be weakly compressive. Using a power spectrum specified by Tu and Marsch, he calculated the expected power spectrum of density fluctuations, if the waves propagate obliquely. At high wavenumbers, the predicted density spectrum at $r = 5r_S$ is 2–3 orders of magnitude larger than the observations reported by Coles & Harmon (1989). Unless the waves really are nearly parallel-propagating, which is difficult to

imagine in a structured corona, we conclude that the direct-launching scenario is not viable. (See HI for a discussion of other objections to direct launching.)

A third proposal for the origin of the ion-cyclotron waves has emerged in recent years, viz., the waves are locally generated by plasma microinstabilities in the corona or transition region (e.g., Viñas *et al.* 2000; Markovskii 2001; Markovskii & Hollweg 2002a; Voitenko & Goossens 2002a, 2003, 2005; Chen & Zhou 2003). The unstable waves usually propagate highly oblique to the background magnetic field. In some studies the instabilities are driven by the currents or gradients associated with low-frequency waves which contain most of the power. But the unstable waves are cyclotron resonant with protons and ions. The net result is a direct transfer of wave energy from low to high frequencies, without a ‘cascade’ through intermediate frequencies.

In other studies the instabilities are driven by electron or proton beams. An example of this type of study (one which we believe to be particularly promising) is the work of Markovskii & Hollweg (2002b, 2004a, 2004b). They assume that microflares near the coronal base intermittently launch bursts of large electron heat flux outward into the corona. Via a Landau resonance with the distorted thermal electron distribution function, the heat flux drives electrostatic or electromagnetic ion-cyclotron waves unstable. These waves are highly oblique to the magnetic field, but with sunward $\omega/k_{\parallel}$ (in the local plasma frame); as discussed above in connection with Fig. 3, these sunward-propagating waves could be crucial to the behavior of O^{+5} . The cyclotron waves give perpendicular proton and ion heating. (Since the generated waves are highly dispersive, the particles diffuse approximately along hyperbolae in velocity space, not the circles used in our previous examples.)

Markovskii *et al.* (2006) have proposed a completely different idea, one which does not involve the generation of waves with high $k_{\parallel}$. They note that a turbulent cascade involves velocity-shears which in turn drive up instabilities with high $k_{\perp}$. These unstable waves have resonances at harmonics of the proton–cyclotron frequency. This model agrees with several observational features of the turbulent spectrum:

1. The theoretical spectrum tends naturally toward k^{-3} in the dissipation range, as is frequently (but not always) observed.
2. As predicted by the model, the observed values of the wavenumber at which the spectrum transitions from the inertial range into the dissipation range correspond to the same dimensionless “velocity shear parameter” under a wide range of conditions.

7. A few promising recent ideas

In section 4, we mentioned that some models were internally inconsistent in that they assumed purely outward-propagating Alfvén waves in the WKB limit. In that case the nonlinear terms which lead to turbulence sum to zero. For turbulence to develop, there must be a mix of inward- and outward-propagating waves, or the waves must be non-WKB. Dmitruk *et al.* (2002) proposed a simple model based on equations which describe the non-WKB propagation of linear Alfvén waves, but with an additional nonlinear term giving a dissipation rate dimensionally similar to the Kolmogorov rate, but with the important difference that there is no dissipation unless both inward- and outward-propagating waves are present. Dmitruk *et al.* (2002) offer some numerical calculations of the volumetric heating rates in the corona; the derived heating rates are

comparable to what is required to drive the fast wind. We believe that this approach contains a significant amount of the physics needed for describing how the Alfvén wave amplitudes evolve as they propagate away from the Sun, and for describing the overall heating in the corona and interplanetary space; at the same time, this approach offers the virtue of simplicity. (See also Matthaeus *et al.* (1999); Dmitruk *et al.* (2001); and Dmitruk & Matthaeus (2003).) Cranmer & van Ballegoijen (2005) have shown that the approach of Dmitruk *et al.* (2002) accounts well for what is known about the evolution of wave amplitudes throughout the corona and interplanetary space. However, there is a note of caution: The Dmitruk *et al.* model presupposes a cascade to high $k_{\perp}$ rather than high $k_{\parallel}$; is there some way to modify it along the lines of Chandran (2005)?

The reader might also wish to consult Isenberg *et al.* (2003); Matthaeus *et al.* (2004); and Breech *et al.* (2005) for examples of calculations of the evolution of turbulent fluctuations in the solar wind far from the Sun, where V_A can be neglected compared to the solar wind flow speed. Unlike Dmitruk *et al.* (2002), these models include the effects of velocity shear and pick-up ions. Velocity shear may be an important aspect of the evolution, perhaps accounting for some of the differences between the evolution of waves/turbulence in fast and slow solar wind flows (Breech *et al.* 2005).

A wave-turbulence phenomenology such as we have just discussed can describe many aspects of the wave evolution and plasma heating in the fast wind. But there are other aspects of the wave evolution which turbulence phenomenology does not capture. It should first be noted that if waves supply the energy for the wind, then they must be very linear ($|\delta \mathbf{B}|/B \ll 1$) in the vicinity of the acceleration region; otherwise there would be more energy supplied than can be accounted for. However, even allowing for turbulent dissipation, the tendency to conserve wave action implies that the waves become nonlinear ($|\delta \mathbf{B}|/B \approx 1$) in interplanetary space, as is observed. But even though the magnetic field direction fluctuates strongly, the magnetic field strength is observed to be nearly constant ($|\delta \mathbf{B}|/B \ll 1$); in other words, the tip of the magnetic field vector moves nearly on a sphere. How this comes about has been explained via a nonlinear wave analysis combined with hybrid simulations (Vasquez & Hollweg 1996a, 1996b). Nonlinear wave theory (Vasquez & Hollweg 1998, 1999a) has also succeeded in explaining why there are many apparently stable rotational discontinuities (essentially sharp-crested Alfvén waves) imbedded within the waves in interplanetary space (Tsurutani *et al.* 1994), even though an isolated rotational discontinuity would be unstable and very short-lived. Moreover, even in high-speed wind, where the Alfvén waves are most ‘pure’, there is a weak admixture of fast waves propagating across the background magnetic field and pressure-balanced structures (or perhaps transversely-propagating slow modes) (Tu & Marsch 1994); this result has also been explained as a natural outcome of the nonlinear evolution of Alfvén waves (Vasquez & Hollweg 1999b). Another result best explained using wave theory concerns the behavior of He^{++} . We have already said that the He^{++} in interplanetary space flows faster than the protons, by about V_A . But V_A is a decreasing function of r in interplanetary space, implying that the He^{++} must slow down. This, too, has best been explained in terms of He^{++} interacting with nonlinear waves (Kaghashvili *et al.* 2003).

Thus far, turbulence theory has not addressed what happens in the dissipation range, where non-MHD kinetic processes are critical. Similarly, nonlinear wave studies have not simultaneously reproduced a turbulence cascade and the details of the dissipation range. The goal would be to calculate how the proton and ion distribution functions

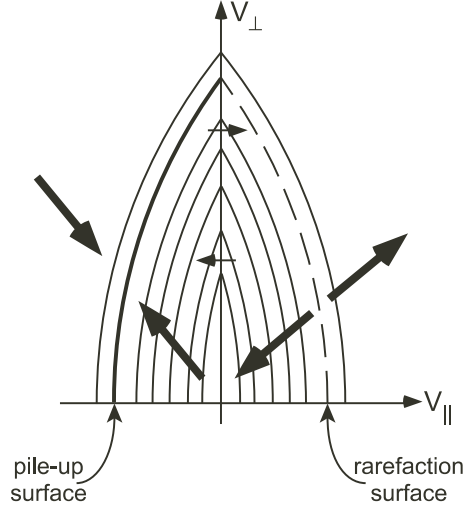


Figure 4. Velocity space diagram illustrating the evolution of the proton distribution in the kinetic shell model, as the plasma moves away from the Sun. Protons interacting with outward (inward) propagating waves diffuse along the arcs in the left (right) half of the diagram. The thick arrows indicate the shell motion, converging on the sunward ($V_{\parallel} < 0$) side and diverging on the antisunward side ($V_{\parallel} > 0$). The thin arrows indicate the particle transport across the $V_{\parallel} = 0$ boundary, determined by the nonresonant forces. The pile-up and rarefaction surfaces are shown as the thick and dashed curves, respectively.

evolve subject to resonant diffusion in velocity space, along with global forces such as gravity, magnetic mirroring, etc. Isenberg *et al.* (2001) and Isenberg (2001b, 2004a, 2004b) have suggested an approximate procedure, which is mostly analytical. They call their approach “the kinetic shell model”, the shells being the 3-dimensional versions of the diffusion arcs in Fig. 2. (See also Galinsky & Shevchenko 2000 for independent work which is closely related to the kinetic shell model.) The essence of this model is shown in Fig. 4, which is similar to Fig. 2. Velocity space ($V_{\parallel}$, $V_{\perp}$) is displayed with several diffusion arcs for protons. The kinetic shell model’s key assumption is that the velocity space diffusion is faster than any other timescale, so that the distribution function is always very nearly uniform along each arc. However, the number of particles can differ from shell to shell. Individual shells can be thought of as moving in response to the global forces: gravity, mirroring, and the charge-separation electric field.

Consider first sunward-moving protons, which we take to have $V_{\parallel} < 0$. Recall that these protons resonate with outward-propagating waves. The leftmost shells, which reach to large values of $V_{\perp}$, experience a net mirror force which overwhelms gravity and the electric field, so those shells are pushed to larger values of $V_{\parallel}$, as indicated by the leftmost arrow. Other shells having $V_{\parallel} < 0$, but reaching to smaller values of $V_{\perp}$, are dominated by the sunward combination of gravity and electric field; those shells are driven to more negative values of $V_{\parallel}$, as indicated by the second arrow from the left. The result is that the sunward-moving protons tend to accumulate near a ‘pile-up surface’ where all forces balance.

At $V_{\parallel} = 0$ the protons are resonant with waves having infinite wavenumber, and zero power. Thus, those protons are unaffected by the waves, and in the kinetic shell model

they move across the $V_{\parallel} = 0$ ‘boundary’ in response to the global forces. If $V_{\perp}$ is large enough, the particles will be pushed toward $V_{\parallel} > 0$, as indicated by the right-pointing arrow; conversely for particles with lower $V_{\perp}$. (Effects such as resonance broadening can also transport protons across $V_{\parallel} = 0$ (e.g., Gary & Saito 2003), but they are not included in the kinetic shell model.)

Particles with $V_{\parallel} > 0$ resonate with sunward-propagating waves, and follow the diffusion arcs sketched in the figure. The two rightmost arrows indicate the shell motions in response to the global forces; shells which reach to large $V_{\perp}$ are pushed to larger values of $V_{\parallel}$, and so on. In this case the opposite of a pile-up surface develops: there will be a ‘rarefaction surface’ on which particles are depleted. Even if there are no sunward-propagating waves initially, in this model they will be generated by those protons at large $V_{\perp}$ which are pushed into the $V_{\parallel} > 0$ region by the mirror force. The resulting distribution is highly unstable to the generation of sunward-propagating waves; the protons appearing at $V_{\parallel} > 0$ will diffuse down and to the right in the figure, losing energy which is made available to the sunward waves.

With dispersion ignored (Isenberg 2001b), rapidly accelerating fast winds were obtained. But when wave dispersion was included (Isenberg 2004a, 2004b), only slow winds were produced. The reason has to do with the tops of the arcs in Fig. 4, which tend to flatten out when the waves are dispersive. Thus the arcs extend to smaller values of $V_{\perp}$, and the net mirror force is reduced. Nonetheless, in the author’s opinion the kinetic shell model has many virtues. Perhaps further developments including the warm plasma dispersion relation and obliquely-propagating waves will save the day.

8. Summary and inferences

We have shown how our thinking about the solar wind has progressed from Parker’s electron-driven wind, through a wave-driven wind, to our current proton-driven wind, driven mainly the magnetic mirror force if the protons have $T_{\perp} \gg T_{\parallel}$. (See Vázquez *et al.* (2003) for a discussion of the importance of thermal anisotropy.) Our current picture also allows for a significant contribution from wave pressure, but this will give a more gradual acceleration acting mainly beyond the sonic point. We have emphasized that the wind is undergoing significant heating in the region where it is rapidly accelerating to supersonic speeds; coronal heating and solar wind acceleration need to be treated together.

The UVCS/SOHO fast solar wind data have shown that coronal heating works mainly on the transverse components $\delta V_{\perp}$ of protons and ions. It is not Joule heating, as is very commonly presumed to be the case in other parts of the corona. Coronal heating also does not seem to be dominated by viscosity or by heat conduction. Do proton and ion heating dominate other parts of the corona, such as the active region loops, as well?

Heating transverse to the magnetic field strongly implicates the ion-cyclotron resonance. But we still do not know the source of the high-frequency resonant waves. We have given some arguments against the proposal that the Sun directly launches these waves. The Sun could well launch the required energy fluxes at long periods, with the high-frequencies generated by a turbulent cascade. The tendency of MHD turbulence to produce mainly high $k_{\perp}$ is a difficulty, but we do not believe it is insurmountable. Finally, some workers have begun to explore the promising idea that the high-frequency

waves may originate locally in the corona from plasma microinstabilities; this scenario seems worth pursuing.

In the foregoing we have made several references to radio studies of the corona. They provide evidence for a turbulent cascade to high $k_{\parallel}$ close to the Sun, and we shall see that they provide evidence for electron heating via Landau damping of long-period fluctuations. They are the only means we have of measuring the all-important magnetic fluctuations in the corona, and they are therefore the only means we have of determining whether long-period MHD waves can provide, via a turbulent cascade, sufficient energy to heat the protons and ions and drive the high-speed wind. (If radio studies were to show that there is insufficient power at long periods, then the turbulence scenario would be dead.) We believe that the solar physics community should pay more attention to radio studies than has generally been the case. The author suggests that spacecraft which will be close to superior conjunction with the Sun should be fitted with linearly polarized transmitters to facilitate observations of the coronal magnetic field via Faraday rotation; we cannot emphasize too strongly that, short of a “Solar Probe”, this is our only hope of measuring coronal magnetic fluctuations. We would even encourage the placement of a simple spacecraft with little more than a linearly polarized transmitter in an orbit which places it in near constant superior conjunction (Ruzmaikin *et al.* 1997).

We also encourage development of an advanced version of the UVCS, especially one which will allow us to get a handle on the behavior of helium in the acceleration region of the solar wind. A complementary instrument with great promise is the ultraviolet imager discussed by Kohl (this volume). And the author hopes to live long enough to see a “Solar Probe”.

We hope that our presentation has made it clear that a complete description of coronal heating and fast wind acceleration will require detailed considerations of particle kinetics; MHD does not tell the whole story. The non-Maxwellian particle distribution functions observed by spacecraft (e.g., Marsch 1991) are trying to tell us something about the physics of heating and acceleration closer to the Sun.

Future studies will need to place more emphasis on wave propagation oblique to the magnetic field. The Sun almost certainly launches highly oblique waves which are subsequently refracted. The microinstabilities considered so far tend to produce highly oblique waves. Oblique waves in a multi-ion plasma show interesting polarization effects, for example at the “crossover frequencies” (e.g., HI; Smith & Brice 1964). There is some evidence (Leamon *et al.* 1998a, 1998b, 1999, 2000; Bale *et al.* 2005) that the dissipation range of the turbulence contains kinetic Alfvén waves, which are essentially highly oblique Alfvén waves (Hollweg 1999) with cross-field wavenumbers of the order of or greater than the inverse ion inertial length. These waves are compressive and therefore subject to Landau damping, especially on the electrons. Voitenko and Goossens (2002b, 2004) have emphasized the possible importance of kinetic Alfvén waves. The latter paper considers what happens to test ions interacting with a large-amplitude kinetic Alfvén wave packet. Surprisingly, under certain circumstances the ions can gain substantial energy even though there is no cyclotron or Landau resonance. This process might be important far from the Sun where the solar wind wave field has large amplitude, but it will not be important close to the Sun where the waves are almost certainly small amplitude with $|\delta \mathbf{B}|/B \ll 1$. On the other hand, Wu & Yang (2007) have recently studied heavy ion energization in a “solitary non-linear [kinetic Alfvén] wavelet”. They find that significant energization can occur even

if the magnetic field variation is weak, as long as the electron density varies strongly within the wavelet.

Future studies will also need to expand consideration of sunward-, as well as outward-propagating waves. The mix of inward and outward waves is essential for the development of turbulence. Moreover, we believe that inward-propagating waves are responsible for the extended heating of O^{+5} .

Finally, we need to say something about electron heating, which has been completely ignored in the foregoing. It is difficult to say much about the electrons because we do not understand how to model their heat conduction under weakly collisional conditions. It is possible that the electrons need no external heating, and that collisional coupling with the protons suffices to maintain their temperatures. It is also possible that the electrons are strongly heated intermittently by reconnection events at the coronal base. And it may be that electrons are heated throughout the corona via kinetic Alfvén waves, which are compressive and have a parallel electric field; since the electron thermal speed is comparable to V_A in the corona, the electrons can be heated via the Landau resonance (e.g., Leamon *et al.* 1999, 2000). From their radio studies close to the Sun, Harmon & Coles (2005) conclude “at least half of the total cascade flux goes into electron heating as cascaded power traverses the broad Landau resonance on the way to the cyclotron resonance”. It has also been suggested that electrons may be heated by nonlinear dynamics at very small scales, perhaps in association with reconnections in the dissipation range of a turbulent cascade (Matthaeus *et al.* 2003) or in association with kinetic Alfvén waves (Cranmer & van Ballegooijen 2003).

Fisk (2003) and Gloeckler *et al.* (2003) have emphasized observational evidence that the fastest (slowest) solar wind originates from the coronal regions with the coolest (hottest) electrons, a result completely opposite to the model of Parker (1958). They explain this result in terms of ‘interchange reconnections’ which convert closed loop-like magnetic field lines into open field lines. Plasma originally on the loops is thereby made available to the solar wind. Space does not permit us to reproduce their arguments, but the net result is that the hotter loops contain more mass than the cooler loops, and with more mass to accelerate, slower flows result. Alternatively, Schwadron and McComas (2003) have suggested that coronal regions with hotter electrons give rise to slower solar wind because those regions conduct more energy down to the transition region where it is lost via radiation, leaving less energy for the kinetic energy of the flow. Cranmer *et al.* (2007) have emphasized the magnitude and shape of the magnetic field expansion, which determines where heat is deposited in the corona. Finally, the electron temperature-speed anticorrelation might be consistent with the idea that it is mainly the protons which are heated (by the mechanisms discussed in this paper) in all regions which give rise to the solar wind. The protons then heat the electrons via collisions. Less dense regions, viz., the coronal holes, have weaker collisional coupling between electrons and protons, and this presumably results in lower electron temperatures coupled with faster flows.

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Numerical Simulations of Kinetic Alfvén Waves to Study Spectral Index in Solar Wind Turbulence and Particle Heating

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Abstract. We present numerical simulations of the modified nonlinear Schrödinger equation satisfied by kinetic Alfvén waves (KAWs) leading to the formation of magnetic filaments at different times. The relevance of these filamentary structures to solar wind turbulence and particle heating has also been pointed out.

Key words. Kinetic Alfvén wave: filamentation—solar wind—spectral index—chaotic structure.

1. Introduction

Alfvén wave turbulence is commonly observed in space plasmas, especially in the solar wind, the magnetosheath and auroral regions. The kinetic Alfvén wave (KAW) is obtained when the magnetohydrodynamic (MHD) Alfvén wave develops a large wavenumber, $k_{\perp}$, transverse to the ambient magnetic field B_0 . The KAW carries nonzero parallel electric and magnetic field perturbations which contribute to plasma energization and particle acceleration. One such means for transferring energy from the large scale to small scale is the transverse collapse leading to the formation of strong magnetic filaments parallel to the ambient field, as asymptotically predicted by the nonlinear Schrödinger equation for the wave envelope (Champeaux *et al.* 1997, 1998; Laveder *et al.* 2001, 2002).

The motivation of this article is to solve the modified nonlinear Schrödinger equation (MNLSE) satisfied by the KAW numerically and study the transient filamentation process arising on account of the coupling between the main KAW and the perturbation.

2. Model equations

The dynamical equation governing the propagation (in the x – z plane) of low frequency, long wavelength, and finite amplitude KAW (magnetized by a uniform ambient magnetic field B_0 along the z direction), for intermediate plasma beta ($m_e/m_i < \beta < 1$; $\beta = 8\pi n_0 T/B_0^2 < m_e/m_i$) is governed by the MNLS equation (Bellan & Stasiewicz 1998; Shukla & Stenflo 1999, 2000; Shukla *et al.* 1999; Shukla & Sharma 2002) in dimensionless form,

$$i \frac{\partial B_y}{\partial t} + i \frac{\partial B_y}{\partial z} + 2i\Gamma \frac{\partial B_y}{\partial x} + \frac{\partial^2 B_y}{\partial x^2} + |B_y|^2 B_y = 0, \quad (1)$$

where $\partial_z B_y \ll k_{0z} B_y$, $m_e(m_i)$ is the mass of the electron (ion), n_0 is the unperturbed plasma number density, $T(=T_e \approx T_i)$ is the plasma temperature. $\Gamma = (v_{te}/v_A)k_{0x}\lambda_e$ is a parameter characterizing the normalized perpendicular wave number in terms of electron's collisionless skin depth, $v_{te}(= \sqrt{T_e/m_e})$ is the electron thermal speed, $v_A(= \sqrt{B_0^2/4\pi n_0 m_i})$ is the Alfvén speed, $k_{0x}(k_{0z})$ is the component of the wave vector perpendicular (parallel) to $\hat{\mathbf{z}}B_0$, and $\lambda_e(= \sqrt{c^2 m_e/4\pi n_0^2})$ is the collisionless electron skin depth. The normalizing values are $t_n = (2\omega/v_A^2 k_{0z}^2)(1 + k_{0x}^2 \lambda_e^2)$, $z_n = 2/k_{0z}$, $x_n = v_{te} \lambda_e/v_A$, and $B_n = [1 - \Delta(1 + \delta)]v_A^2 k_{0z}^2/16\pi n_0 T_e \omega^2]^{-1/2}$ where ω is the Alfvén wave frequency, $\Delta = \omega^2/\omega_{ci}^2$, $\delta = m_e k_{0x}^2/m_i k_{0z}^2$, and $\omega_{ci}(= cB_0/m_i c)$ is the ion gyrofrequency.

3. Numerical simulations

We have carried out 2D pseudospectral simulation of equation (1) numerically in a $(2\pi/\alpha_x) \times (2\pi/\alpha_z)$ periodic domain with $\alpha_x, \alpha_z = 1$ and $(128)^2$ grid points with initial condition of simulation

$$B_y(x, z, 0) = |B_{y0}|(1 + 0.1 \cos(\alpha_x x))(1 + 0.1 \cos(\alpha_z z)), \quad (2)$$

where B_{y0} is the amplitude of the homogenous pump KAW. The finite difference with predictor–corrector method was used for the integration of time with a small time step of $dt = 5 \times 10^{-4}$. The results of filament formation of KAW for various times and keeping $\Gamma(= 0.01)$ fixed are presented below.

The time evolution of the intensity of the transverse magnetic field is exemplified in Fig. 1(a–d) by means of snapshots at four instants of time ($t = 1.5, 2, 2.5, 4.5$). It is evident from the figure that at early times more intense filaments are formed and with the advancement of time, these become chaotic structures. Once the chaotic structure is formed, the filaments become more or less chaotic at later times.

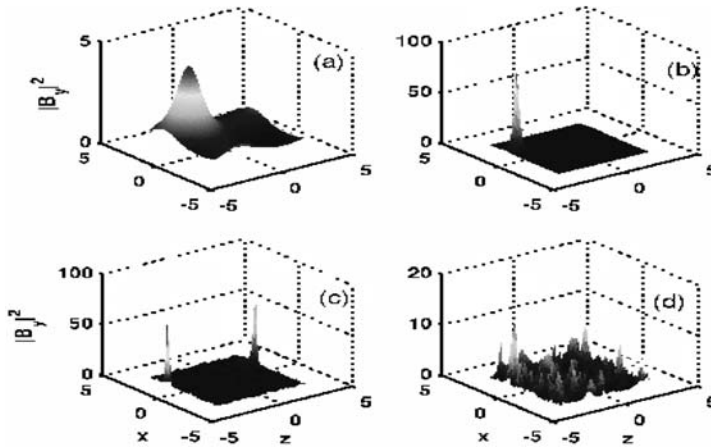


Figure 1. The magnetic-field intensity profiles of KAW. (a) $t = 1.5$, (b) $t = 2$, (c) $t = 2.5$, and (d) $t = 4.5$.

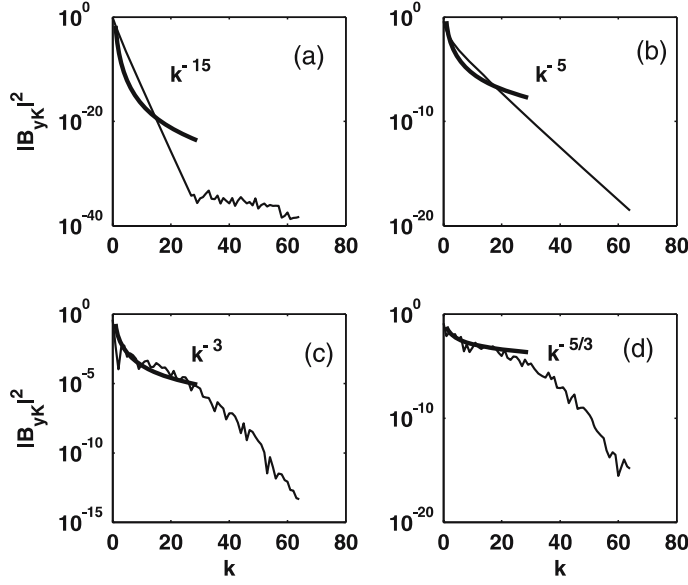


Figure 2. Variation of $|B_{yk}|^2$ against k_z at $k_x = 0$ for KAW. (a) $t = 1.5$, (b) $t = 2$, (c) $t = 2.5$, and (d) $t = 4.5$. The thick line curve indicates the spectral index scaling.

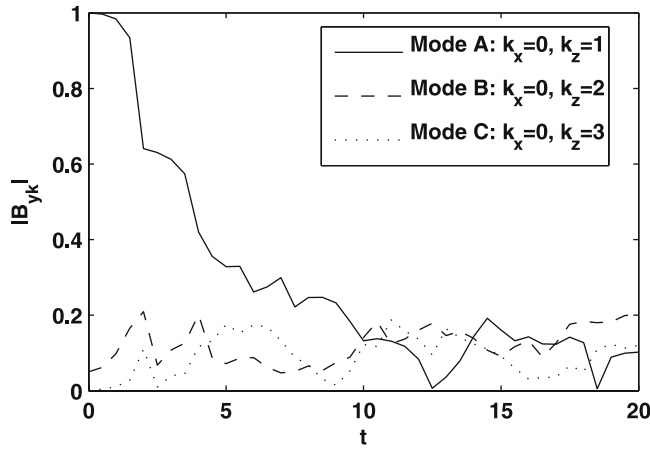


Figure 3. Time evolution of $|B_{yk}|$ of three Fourier modes for KAW.

Figure 2(a–d) shows the variation of $|B_k|^2$ against k_z for a fixed value of $k_x = 0$ at different times. It is evident from the wavenumber spectrum that the spectral index is steeper than the Kolmogorov $k^{-5/3}$ scaling at early times. But as time advances, it approaches Kolmogorov $k^{-5/3}$ scaling. This is in agreement with the power spectra for various time series found by Goldstein *et al.* (1999) that for late time and large distances, the spectral index approached Kolmogorov $k^{-5/3}$ scaling.

Figure 3 shows the time evolution of the amplitudes of several wave modes and represents that the nonlinear growth of the mode A occurs at $t = 12.5$. Since the $k^{-5/3}$

scaling has already formed at $t = 4.5$, the growth of the mode A at $t = 12.5$ occurred under the turbulent state of waves. Modes B and C grow first and decay showing the oscillatory evolution.

For application purposes, the typical values of several solar wind parameters as measured by Helios 2 at 1 AU are $B_0 \approx 6$ nT, $n_0 \approx 3$ cm⁻³, $T \approx 10$ eV; then $\beta \approx 0.335$, $v_A \approx 7.7 \times 10^6$ cm/s, $v_{te} \approx 1.3 \times 10^8$ cm/s, $\omega_{ci} \approx 0.6$ Hz, and $\rho_s \approx 5.3 \times 10^6$ cm. For these typical parameters at $\omega/\omega_{ci} = 0.02$ and $k_x \rho_s = 0.01$, the magnitudes of the fluctuations are $|\delta B_{y0}|/B_0 = 0.08$ and $\Gamma = 0.01$. In the solar wind region and over a wide range of heliospheric distances, the quantity $|\delta B_{y0}|/B_0$ tends to be small ($\approx 5\%$) (Matthaeus & Goldstein 1982; Roberts *et al.* 1987). The observed filamentary structures of Fig. 1 have a characteristic scale size of the order of the ion gyroradius ρ_s .

4. Discussion and conclusion

We have investigated the transient filamentation instabilities of large-amplitude, KAWs propagating along the background magnetic field. These filaments may act as the source for parametric instabilities of the Alfvén waves and may act as a source of decay waves as well as a source of further collapse. The transverse collapse (filamentation) of KAW produces small scale generation which allows dissipation processes like ion-cyclotron resonance or Landau damping to act, leading to the heating of the plasma. Self-consistent models can be developed based on the kinetic theory to estimate the plasma heating. The wave number spectrum can be used to estimate the additional heating by calculating the velocity space diffusion coefficient in Fokker-Planck equation.

Acknowledgement

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Damping of Slow Magnetoacoustic Waves in an Inhomogeneous Coronal Plasma

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Abstract. We study the propagation and dissipation of slow magnetoacoustic waves in an inhomogeneous viscous coronal loop plasma permeated by uniform magnetic field. Only viscosity and thermal conductivity are taken into account as dissipative processes in the coronal loop. The damping length of slow-mode waves exhibit varying behaviour depending upon the physical parameters of the loop in an active region AR8270 observed by TRACE. The wave energy flux associated with slow magnetoacoustic waves turns out to be of the order of $10^6 \text{ erg cm}^{-2} \text{ s}^{-1}$ which is high enough to replace the energy lost through optically thin coronal emission and the thermal conduction below to the transition region. It is also found that only those slow-mode waves which have periods more than 240 s provide the required heating rate to balance the energy losses in the solar corona. Our calculated wave periods for slow-mode waves nearly match with the oscillation periods of loop observed by TRACE.

Key words. Sun: corona—active region—coronal loops—TRACE—MHD waves—heating.

1. Introduction

The temperature of the solar corona is of the order of 10^6 – 10^7 K and it is much higher than the underlying photospheric temperature (6000 K), indicating the presence of heating mechanisms. Detailed reviews on this topic are in Narain & Ulmschneider (1990, 1996), Browning (1991), and Zirker (1993). MHD waves play an important role in solar and astrophysical contexts. An important class of coronal heating mechanism is the magnetohydrodynamic (MHD) wave heating mechanism, which relies on the dissipation of wave energy to provide the necessary heating (Osterbrock 1961; Gorden & Hollweg 1983; Kumar *et al.* 2006; Kumar & Kumar 2006). MHD waves generated by footpoint motions of coronal loops are capable to carry the sufficient amount of energy, and their dissipation via viscosity, thermal conductivity, resistivity, and radiation heats the coronal magnetic loops. Aschwanden (2004a) pointed out that MHD oscillations and waves with acoustic phase speeds carry insufficient energy to contribute significantly to coronal heating, whereas MHD oscillations and waves with Alfvénic phase speed carry an energy flux that is comparable to the coronal losses due to radiation and thermal conduction, and thus potentially contribute to the heating

of active region loops or postflare loops. Coronal MHD waves, oscillations and solar coronal heating theories have been discussed by several researchers (Aschwanden 2004b; Nakariakov 2003). In this paper, we will study the dissipation of slow MHD waves in an inhomogeneous, compressible, and low- β coronal loop plasma through viscosity and thermal conduction. The paper is organized as follows. In section 2, we present a summary of coronal loop observations taken by TRACE, and the estimations of total energy losses in the coronal loops. The description of the physical model, the governing equations, and general dispersion relation are presented in section 3. In section 4, dispersion relation is solved numerically and the results for different coronal loop parameters are shown. Finally, conclusions are drawn in section 5.

2. Summary of observations and energy losses in coronal loops

Aschwanden *et al.* (1999) have analyzed a sequence of TRACE 171 Å and 195 Å images of an active region AR 8270 on 14 July 1998 during which a flare (GOES M4.6) was reported with a start time around 12:55 UT in the temperature range of 1.0–1.5 MK. The oscillating loops show evidence of strong damping, the amplitude of the oscillations decreases by more than 50% in several oscillation periods. The geometric and physical parameters of transverse oscillations in 26 coronal loops have also been measured by Aschwanden *et al.* (2002) in the 171 Å and 195 Å wavelength bands observed by TRACE. The average values of parameter for the oscillating loops reported by Aschwanden *et al.* (1999, 2002) observed by TRACE are given in Table 1.

For the optically thin part of the atmosphere, such as the transition region and the corona, the radiative losses per unit volume are (Priest 1982):

$$Q_{\text{rad}} = 10^{-18.66} T^{-1/2} n_e^2, \quad (1)$$

the conductive losses based on the assumption made by Krall (1977) can easily be derived,

$$Q_{\text{cond}} = \frac{5.14 \times 10^{-7} T^{7/2}}{2L^2}, \quad (2)$$

where L is the loop half length. The sum of these two loss terms (radiative loss and conductive loss terms) gives the minimum heating rate necessary to sustain the coronal

Table 1. Loop parameters reported by Aschwanden *et al.* (1999, 2002) observed by TRACE.

Parameter	TRACE	
Loop length	110 ± 53 Mm	130 ± 30 Mm
Loop width	8.7 ± 2.8 Mm	7.2 ± 0.80 Mm
Oscillation period	321 ± 140 s	276 ± 25 s
Decay time	580 ± 385 s	
Electron density	$(6.0 \pm 3.3)10^8 \text{ cm}^{-3}$	$(2.0 \pm 0.6)10^8 \text{ cm}^{-3}$
Temperature	$(1.0\text{--}1.5)10^6 \text{ K}$	$(1.0\text{--}1.5)10^6 \text{ K}$
Magnetic field	3.0–30.0 G	$\sim 20.0 \text{ G}$
Reference	Aschwanden <i>et al.</i> (2002)	Aschwanden <i>et al.</i> (1999)

loops at a temperature of the order of 2 MK. Corresponding to the loop parameters discussed in Table 1, the minimum required damping rate for slow-mode waves to sustain the coronal loops at such high temperatures can be determined by equating the volumetric wave heating rate to the sum of the radiative and the conductive losses as $\nu U = Q_{\text{rad}} + Q_{\text{cond}}$, where ν is the wave energy heating rate and is twice the wave-amplitude damping rate, $\text{Im}(\omega)$, U is the mean energy density of the wave and can be determined as the product of the equilibrium mass density ρ_0 and the square of the rms velocity amplitude.

3. Physical model and dispersion relation

We consider the coronal loop as a static, straight, gravitation-less, low- β plasma slab with half-length L along the Z -axis of the cartesian coordinate system. The plasma is permeated by a uniform straight magnetic field ($\mathbf{B} = B_0 \hat{e}_z$) and the density is inhomogeneous along z -direction. We have assumed that there is no background flow $\mathbf{V}_0 = 0$, and viscosity and thermal conduction are considered as dissipative mechanisms. The inhomogeneity of the plasma is introduced by a continuously varying density

$$\rho_0(z) = \rho_c \exp\left(\frac{-z}{L}\right), \quad (3)$$

where ρ_c is the coronal mass density and z is the height from coronal base. Therefore, $T_0(z) \sim p_0/\rho_0$, implies that the plasma temperature T_0 attains higher value at the loop top.

The basic one-dimensional MHD equations governing the plasma motion in coronal loops are the equation of continuity, equation of momentum, energy equation, and equation of state, used by De Moortel & Hood (2003).

When we linearize the set of equations (1)–(4) of De Moortel & Hood (2003) under first-order approximation and assuming that all the variables are of the form $f = A \exp i(kz - \omega t)$, we obtain the following dispersion relation for the wavenumber k as a function of frequency ω :

$$Lk^4 + Mk^3 + Nk^2 - \omega^3 = 0, \quad (4)$$

where

$$L = \frac{4\eta_0}{3}(\gamma - 1)\kappa_p \frac{T_0}{p_0\rho_0}\omega + i(\gamma - 1)\kappa_p \frac{T_0}{\rho_0},$$

$$M = (\gamma - 1)\kappa_p \frac{T_0}{\rho_0}k_\rho,$$

$$N = \omega C_S^2 - \frac{4i\eta_0}{3\rho_0}\omega^2 - i(\gamma - 1)\kappa_p \frac{T_0}{p_0}\omega^2.$$

Here all units are considered in cgs. The dispersion relation (4) is a fourth order polynomial in wavenumber k and, it has been solved numerically for the different loop parameters given in Table 1. The numerical solution of equation (4) gives the complex k 's. The damping length of the slow-mode waves is given by the inverse of k_i , i.e.,

the imaginary part of the wavevector k . The mechanical energy flux density of the magnetoacoustic wave, F , given by Pekünlü *et al.* (2001) is

$$F = \rho \langle \delta v^2 \rangle \frac{\delta \omega}{\delta k}, \quad (5)$$

where $\delta \omega / \delta k$ is the group velocity of the wave which is determined from the dispersion relation (4), $\langle \delta v^2 \rangle$ is the square of the rms velocity amplitude.

4. Results and discussion

Figure 1 shows the variation of damping length of the slow-mode waves as a function of height from the coronal base. The damping length is the distance over which the amplitude of the wave drops to $1/e$ times of its original amplitude. We found that the short period waves lie near the base of the corona and an increase in the wave period as a result increases the damping length of the waves. Figure 1 depicts that the damping length of the wave increases very slowly with the height from the coronal base until it reaches up to 9.0×10^8 cm and thereafter the damping length of the wave sharply increases and becomes maximum around $z = 3.0 \times 10^9$ cm (height from the coronal base). The maxima of the curves become more prominent as the wave period increases and shifts towards the right corresponding to the long period waves. A decrease in the damping length is observed for a particular coronal base height range 4.0×10^9 cm to

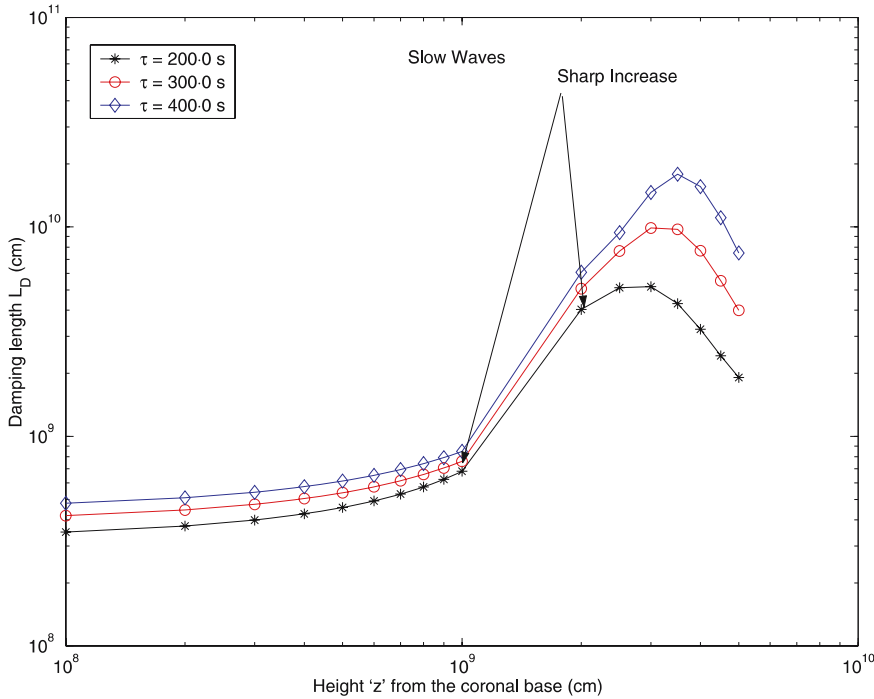


Figure 1. The variation of damping length with height for slow waves with parameters $T_c = 1.0 \times 10^6$ K, $L = 1.0 \times 10^{10}$ cm, and $\rho_c = 1.67 \times 10^{-15}$ gm cm $^{-3}$.

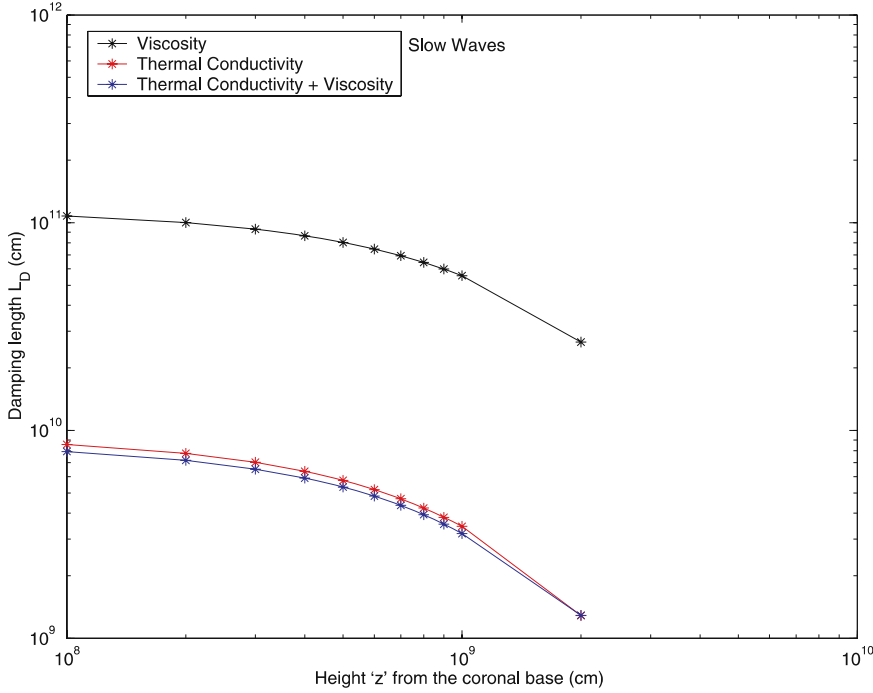


Figure 2. The variation of damping length with height for slow waves with parameters $T_c = 1.0 \times 10^6$ K, $L = 1.0 \times 10^{10}$ cm, $\tau = 200$ s, and $\rho_c = 1.67 \times 10^{-15}$ gm cm $^{-3}$.

6.0×10^9 cm, meaning that mechanical energy is deposited faster in the loop at around (40,000 km). If we consider solar coronal loop with loop half length more than that considered in Fig. 1, then the maxima of the curves shift towards the higher values of z than that of Fig. 1. Figure 2 shows the variation of damping length with the height from coronal base in the coronal loop (Table 1) when different dissipative mechanisms such as viscosity and thermal conductivity are taken into account simultaneously and separately. If we consider the propagating wave in the coronal loop damped through viscosity alone, we find larger damping length in comparison with the loop half length for all possible coronal base heights. Therefore, the dissipation of mechanical energy of the waves via viscosity cannot balance the energy losses in coronal loops. However, if we consider the thermal conductivity alone as a dissipative mechanism in the coronal loop, we find smaller damping length in comparison with the loop half length for all possible heights from the coronal base. It means that the thermal conductivity is more efficient dissipation mechanism than the viscosity within the coronal loops in order to balance the energy losses.

5. Conclusions

Our investigation of the slow magnetoacoustic wave dissipation in the coronal loops leads us to conclude that slow magnetoacoustic waves carry sufficient energy flux to heat the solar corona. If we consider the non-thermal broadening of the coronal lines

as claimed by Saba & Strong (1991), the wave energy flux density turns out to be of the order of $10^6 \text{ erg cm}^{-2} \text{ s}^{-1}$, which is high enough to replace the energy losses via radiation and conduction below to the transition region and chromosphere. We found that only those slow-mode waves which have periods more than 240 s may provide the required heating rate to balance the energy losses in the solar corona. Our calculated wave periods for slow-mode waves in the presence of viscosity and thermal conductivity are consistent with the oscillation periods of loops observed by TRACE. We cannot claim that the wave dissipation is the only effective process to heat the coronal loops because there are many other agents such as magnetic reconnection, current dissipation, phase mixing, resonant absorption, and so on working with different efficiencies in different regions of the solar corona.

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H α Intensity Oscillations in Large Flares

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Abstract. We reinvestigate the problem of H α intensity oscillations in large flares, particularly those classified as X-class flares. We have used high spatial and temporal resolution digital observations obtained from Udaipur Solar Observatory during the period 1998–2006 and selected several events. Normalized Lomb–Scargle periodogram method for spectral analysis was used to study the oscillatory power in quiet and active chromospheric locations, including the flare ribbons.

Key words. Sun: oscillations, chromosphere, flares.

1. Introduction

H α intensity oscillation was detected many decades ago by several workers, e.g., Jensen & Orral (1963) and Elliott (1969) who reported chromospheric oscillations with a period of 120–130s. More recently, Jain & Tripathy (1998) and Wang *et al.* (2000) amongst others showed the influence of flares on the chromospheric oscillations. Considering a variety of contradictory results obtained by earlier workers, we reinvestigate the problem of oscillations in the quiet, active and flaring locations in the chromosphere.

2. The observational data

The H α data of solar flares used for this study were obtained by the 6-inch aperture Spar telescope of Udaipur Solar Observatory in the observational period 04:00–12:00 UT of USO time window. The data have spatial and temporal resolution of the order of 2–3 arcsec and 3s–60s in different datasets. The selection of data was carried out essentially with two criteria: (i) energetic flares belonging to X-class, and (ii) flares occurring in the location between $\pm 50^\circ$ in longitude and $\pm 20^\circ$ in latitude, so that the projection effects may be neglected (Venkatakrishnan & Allen Gary 1989).

3. Data analysis

From the large set of available images, we selected the best images taken at good seeing conditions. Standard image processing steps were then applied on the selected

images. The effects of guiding and image motion were reduced by registering the images. To remove effects of varying observing conditions during the observations we normalized the images.

The relative intensities at quiet (C), plage (P) and flare ribbon (F) points were calculated by taking 8×8 matrix in each frame. These locations were carefully selected using movies of the respective events. The tracking errors and atmospheric seeing effects do not allow use of smaller windows, while larger windows would tend to dilute the net intensity. The power spectrum of these relative intensities were estimated using Lomb (1976) method for data sampled at uneven intervals. Jian *et al.* (2007) have recently used this method for spectral analysis of VLBI observational data, and investigated the problem of appearance of spurious signal in the data with unequally spaced sampling.

4. Results and conclusion

The $H\alpha$ filtergrams for three of the selected flares observed on 28 November 98 (X3.3/3N), 10 April 01 (X2.3/3B) and 26 October 03 (X1.2/1N) were analyzed. $H\alpha$ intensity flare images at maximum phase for the 10 April 01 event is shown in Fig. 1(a). The three positions corresponding to the flare ribbon, plage region and quiet chromospheric region are denoted by F, P and C in the $H\alpha$ filtergram over which we obtained relative intensity profiles as shown in Fig. 1(b), alongwith their respective power spectra in Fig. 2(a–c). The power spectrum of the 10 April 01 flare shows periodicity at 3.3 mHz in quiet chromosphere and in the plage. The peak at 3.3 mHz corresponds to 5-min oscillation in the chromosphere. The power spectrum for the 26 October 03 and 28 November 98 events show dominant peaks at 3.3 mHz indicating the existence of 5-min oscillation in the quiet chromosphere, with other significant peaks at 2.5 and 4.6 mHz. The significant peaks at the plage locations were also found at 3.3 mHz. We found that the amplitudes of some peaks, which were prominent in the quiet chromosphere, reduced in the flare. The other significant peaks at 1 mHz in quiet, plage and flare locations correspond to a spurious large period oscillation perhaps appearing as the artifact of Lomb's method (Jian *et al.* 2007), and is not of interest.

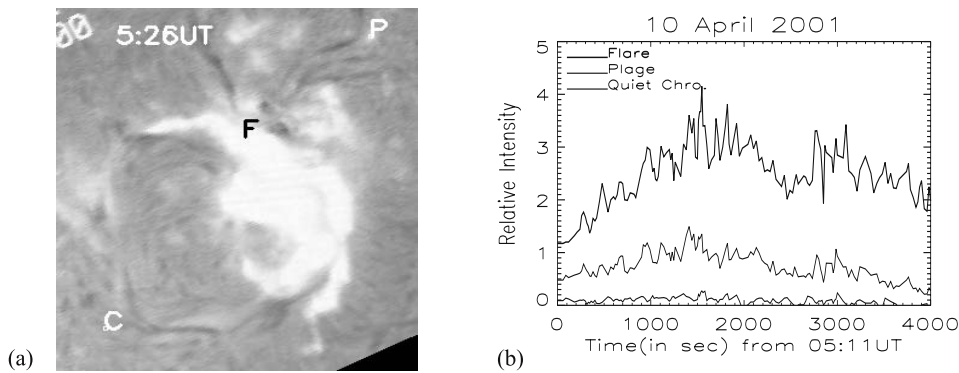


Figure 1. (a) $H\alpha$ intensity flare images at maximum phase – 10 April 01 event. The points F, P and C denote flare ribbon, plage and quiet chromospheric positions over which we obtained relative intensity profiles. (b) $H\alpha$ intensity profiles at three positions F, P and C.

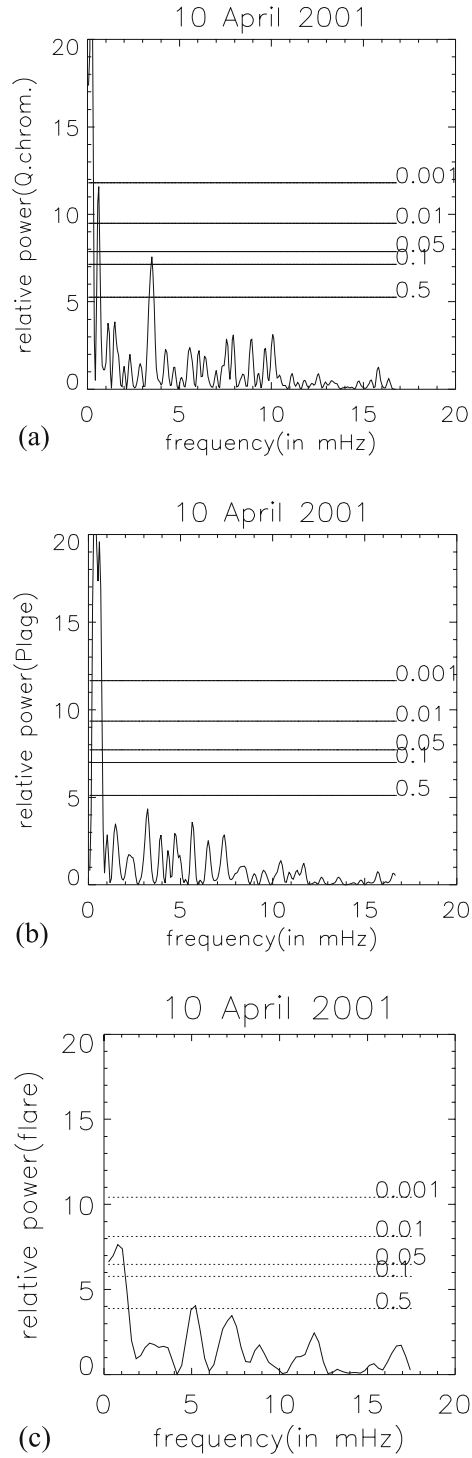


Figure 2. Power spectra corresponding to time profiles at (a) quiet location (reference) ‘C’, (b) plage ‘P’, and (c) flare ‘F’ marked in Fig. 1(a).

In summary, we arrived at the following inferences from H α intensity measurement of these events. The results obtained from the analysis of the six flares are still not conclusive enough and do not consistently provide the 3- and 5-minute oscillations at all the locations observed in H α . However, there are varying levels of detection of these oscillations in individual events, viz.,

- 3- and 5-minute oscillations are present in quiet chromosphere in all the cases alongwith some additional periodicities as reported by earlier researchers.
- These oscillations are modified in the flare locations as found in some cases.
- Nearly the same periodicities are found in plage as in the quiet chromosphere but with reduced amplitude.

We plan to enhance our sample of flares by including M-class flares of significant importance in optical class, such as 1B and larger, in order to increase the statistical significance of the results obtained in this study.

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Role of Magnetic Carpet in Coronal Heating

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Abstract. One of the fundamental questions in solar physics is how the solar corona maintains its high temperature of several million Kelvin above photosphere with a temperature of 6000 K.

Observations show that solar coronal heating problem is highly complex with many different facts. It is likely that different heating mechanisms are at work in the solar corona. The separate kinds of coronal loops may also be heated by different mechanisms.

Using data from instruments onboard the Solar and Heliospheric Observatory (SOHO) and from the more recent Transition Region and Coronal Explorer (TRACE) scientists have identified small regions of mixed polarity, termed magnetic carpet contributing to solar activity on a short time scale.

Magnetic loops of all sizes rise 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 coronal heating. There is enough energy coming up from the loops of the “magnetic carpet” to heat the corona to its known temperature.

Key words. Sun—corona—magnetic field—magnetic carpet—reconnection.

1. Introduction

How the nature of the processes that heat the corona, maintain it at this high temperature and accelerate the solar wind is a great solar mystery. Usually temperatures fall as we move away from a heat source. This is true in the Sun’s interior right up to visible surface. Then, over a relatively small distance, the temperature suddenly rises to extremely high values. This high temperature of corona requires a permanent heating mechanism. There is no definite answer to this question yet, but there are some questions, which can help us to follow this track.

- Where does the energy come from?
- How does the energy get up into the corona?
- How is the energy released in the corona?
- How can we model this heating using physics, mathematics and computer?

There are some clues to help us know how the temperature changes as we go higher in solar atmosphere.

- All energy from the Sun is produced in the solar core – far below the solar atmosphere.
- The laws of thermodynamics suggest that heat energy from a gas of a high given temperature cannot normally be used to heat another gas to a higher temperature.

From the above discussions it is clear that the temperature of corona requires energy by non-thermal processes because simple laws of physics prevent heat from flowing directly from solar surface or photosphere, at about 6000 Kelvin to much hotter corona at about 1 million to 2 million Kelvin.

Two classes of models have been proposed namely magnetic waves (Goossens 1991; Roberts 1991) and magnetic reconnection (Priest *et al.* 2000) to explain the coronal heating.

2. Recent space observations

Using data from instruments onboard the Solar and Heliospheric Observatory (SOHO) (Kohl *et al.* 1997) and from the more recent Transition Region And Coronal Explorer (TRACE) (Nakariakov *et al.* 1999), solar physicists have identified small patches of magnetic field covering the entire surface of the Sun. Contrary to the bright, large magnetic field loops which are linked to the “active regions” during periods of solar maxima, these patches seem to appear and disappear randomly in time scales of the order of 40 hours. Scientists now think that this magnetic carpet is probably a source of the corona’s heat.

According to recent observations, the ‘magnetic carpet’ shows a very dynamical evolution: photospheric magnetic fields constantly move around, interact with each other, dissipate and emerge at very short periods of time.

3. How is reconnection working in corona?

Magnetic reconnection between magnetic fields of opposite polarity may change the topology of the field and release magnetic energy. The reconnection process will also result in the dissipation of electric currents which will transform electric energy into heat.

Understanding how the magnetic field interacts with plasma is the key to many solar phenomena including coronal heating.

Reconnection is a fundamental process in a plasma. It

- changes the topology,
- converts magnetic energy to heat/K.E,
- accelerates fast particles.

The heating of the corona is linked to the interaction of the magnetic field lines radiating out of the small patches mentioned above. Because the laws of electromagnetism prohibit the intersection of two magnetic field lines, every time magnetic field lines come close to crossing they are “rearranged”, and this magnetic reconnection continuously heats the solar corona. It is a fairly inefficient source of energy, but

the sheer number of these small magnetic patches on the surface of the Sun makes the process a viable solution to the 50-year old problem of what heats the solar corona.

Over the years several different mechanisms of heating corona have been proposed using Reconnection Theory as listed below:

- (i) Drive Simple Reconnection (Priest & Forbes 1986; Priest *et al.* 2000)
 - At nullpoint by photospheric motion
 - X-ray bright point supported by TRACE
- (ii) Binary Reconnection (Parker 1979, 1994)
 - Relative motion of pairs of sources in the solar surface
 - Suppose unbalanced and connected.
- (iii) Separator Reconnection (Longcope & Cowley 1996; Longcope & Silva 1998)
 - High order interaction between several sources
 - Initially unconnected
- (iv) Braiding (Parker 1979, 1994)
 - Parker's Model
 - Initial B uniform/motions braiding
 - In this paper we are discussing the Tectonic model of Coronal Heating.

4. Coronal tectonics

Reconnection can be driven at many of the separatrix that separate from each other all the tiny narrow elementary coronal loops whose foot points are the intense flux tubes (Priest *et al.* 2000). The driving will be by the lateral relative motion, the fragmentation and the cancellation of the minute magnetic fragments in the photosphere. The formation and dissipation of current sheets on separatrix surfaces are referred as “flux-tube tectonics”. The behaviour of majority of the photospheric magnetic flux in the quiet Sun is that it emerges as ephemeral regions, which split into network elements. Each network consisting of about 10 unresolved intense flux tubes, and that the flux then migrates to the boundaries of supergranule cells and moves along them, fragmenting, merging and canceling.

4.1 Dissipation of current sheets and rate of heating

The shearing motions of flux elements at the photosphere cause current sheets to grow at the separatrices of the magnetic cells in the corona. These current layers either are very thin (two-dimensional case) or grow straight away as singularities (three dimensional case). Therefore, even ohmic dissipation will soon become very active, possibly enhanced by anomalous resistivity or reconnection in these sheets (Priest & Forbes 1986).

The heating power delivered by the coronal segment of a loop $2L$ and width of the order of the transverse size (2ω) of magnetic cells.

$$H_{\text{loop}} = \frac{4L\dot{u}L^2}{\ddot{a}\sigma}$$

where σ is the effective electrical conductivity of coronal plasma, I is surface current at the separatrix, and δ is sheet thickness in coronal region.

In the carpet region, at a distance r from the flux element, the width of the current channel scales as r/ω , and the contribution to the heating of the region that carries the surface current I from the loop to the flux source element scales as

$$H_{\text{carpet}} = \int_0^\omega 2\pi r r dr \frac{\delta r}{\omega} \frac{I}{\sigma} \left[\frac{I}{\delta(r/\omega)} \right]^2 = \frac{8\pi\omega^2 I^2}{\delta\sigma}.$$

5. Conclusions

The spreading with height above photosphere of the magnetic of the corona which is rooted in the solar surface in myriads of tiny magnetic fragments, intense flux tube of field strength $B = 1200$ G and radius 100 km is likely to have profound implications for the mechanism of coronal heating. Theoretical modeling and the latest space-based observations suggest that unresolved observations of coronal loops should exhibit enhanced heating near their feet in the carpet, while the upper parts of large-scale loops should be heated rather uniformly but less strongly.

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Multi-parametric Effect of Solar Activity on Cosmic Rays

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Abstract. The long-term modulation of cosmic ray intensity (CRI) by different solar activity (SA) parameters and an inverse correlation between individual SA parameter and CRI is well known. Earlier, it has been suggested that the concept of multi-parametric modulation of CRI may play an important role in the study of long-term modulation of CRI. In the present study, we have tried to investigate the combined effect of a set of two SA parameters in the long-term modulation of CRI. For this purpose, we have used a new statistical technique called “Running multiple correlation method”, based on the “Running cross correlation method”. The running multiple correlation functions among different sets of two SA parameters (e.g., sunspot numbers and solar flux, sunspot numbers and coronal index, sunspot numbers and grouped solar flares, etc.) and CRI have been correlated separately. It is found that the strength of multiple correlation (among two SA parameters and CRI) and cross correlation (between individual SA parameter and CRI) is almost similar throughout the period of investigation (1955–2005). It is also found that the multiple correlations among various SA parameters and CRI is stronger during ascending and descending phases of the solar cycles and it becomes weaker during maxima and minima of the solar cycles, which is in accordance with the linear relationship between SA parameters and CRI. The values of multiple correlation functions among different sets of SA parameters and CRI fall well within the 95% confidence interval. In the view of odd–even hypothesis of solar cycles, the strange behaviour of present cycle 23 (odd cycle), as this is characterized by many peculiarities with double peaks and many quiet periods (Gnevyshev gaps) interrupted the solar activity (for example April 2001, October–November 2003 and January 2005), leads us to speculate that the solar cycle 24 (even cycle) might be of exceptional nature.

Key words. Sun—solar parameters—cosmic ray modulation—running multiple correlation method.

1. Introduction

The history of cosmic ray (CR) modulation studies is as long as the history of regular CR observations. Initially, these studies were strictly tied to ground-based neutron monitors. Even now in the epoch of Voyager and Ulysses satellite missions,

when successful CR experiments have been performed into the distant heliosphere, the neutron monitor is still the best detector to investigate CR modulation. Long-term modulation of cosmic rays can be studied by using the monthly data (averages) of global network of neutron monitor stations having different geomagnetic cut-off rigidities. Neutron monitors are most sensitive to cosmic rays in the energy range 0.5–20 GeV, which coincides with the maximum energy response and effective solar modulation. It has been known for a long time that the intensity as well as the energy spectrum of galactic cosmic rays is modulated by solar activity (SA). The details of the CR modulation and variation of time-lag factor are still a matter of great interest. Among various solar activity parameters, the sunspot number (SSN) has been considered as a primary indicator to define the level of solar activity, which in general follows an 11-year periodicity. Therefore, the sunspot numbers are used as a representative solar index for the study of cosmic ray modulation and solar-terrestrial relationship (Dorman & Dorman 1967; Pomerantz & Duggal 1971; Rao 1972; Webber & Lockwood 1988; Ahluwalia 1998). Later on, in addition to SSN, a variety of solar indices (solar flares, solar flux, coronal index, etc.) have been used as a proxy index to represent the solar activity (Kane 2005; Gupta *et al.* 2005, 2006).

Galactic cosmic rays in the energy range from several hundreds MeV to tens of GeV are subjected to heliospheric modulation, under the influence of solar output and its variation. The heliospheric modulation of cosmic ray intensity and spectrum are associated with 11-year solar activity cycle. The charge/polarity dependence of drift mechanism is clearly observed in cosmic ray modulation in terms of 22-year solar magnetic cycle, showing different shapes of cosmic ray maxima in the alternate solar cycles. Long-term cosmic ray modulation in the high-energy range is studied using the monthly mean data of global network of cosmic ray neutron monitoring stations having different geomagnetic cut-off rigidities. Neutron monitors are most sensitive to cosmic rays in the energy range 0.5–20 GeV, which coincides with maximum energy response for effective solar modulation. Though, the anti-correlation between solar activity and flux of galactic cosmic rays reaching the earth is a well established fact, the degree of anti-correlation is found to vary during the different phases of solar cycles (Dorman & Dorman 1967; Pomerantz & Duggal 1971; Rao 1972; Nagashima & Morishita 1979; Mavoromichalaki & Petropoulos 1984; Webber & Lockwood 1988; Nymmik & Suslov 1995; Storini *et al.* 1995; Ahluwalia & Wilson 1996; Dorman *et al.* 2001; Usoskin *et al.* 2002).

In fact, it is unreasonable to expect a single index of solar activity or some heliospheric parameter to affect the cosmic ray variation. The observed CR variations are integrated phenomena by their nature and it is expected that many factors will have an influence in the modulation process of cosmic rays. Hence, a more realistic model of long-term modulation should combine several solar-heliospheric indices (Belov 2000). In the present study, considering the above idea, we have tried to investigate the effect of a set of solar-heliospheric parameters in the modulation of cosmic rays. For such a study, we have performed a detailed correlative analysis between CRI and a different set of SA parameters, using “Running multiple correlation method”. The results obtained by this method have also been compared with the results of correlative analysis between CRI and individual SA parameters.

2. Data and method of analysis

In the earlier studies of cosmic ray modulation a variety of solar-heliospheric parameters have been used to study the linear relationship between CRI and these parameters. In general, sunspot numbers (SSN), solar flux (SF), coronal index (CI), solar flare index (SFI), grouped solar flares (GSF), tilt angle of heliospheric neutral current sheet (TA), etc., have been used (Gupta *et al.* 2006). The results of different SA parameters have also been compared (Mishra *et al.* 2006).

In the present study, we have used the monthly mean values of SSN, SF, CI, GSF and TA as solar-heliospheric parameters for the cycles 19–23. The CRI data has been taken from the mid latitude neutron monitor station located at Kiel. To observe the effect of multiple parameters on CR modulation, we have used a new statistical technique called “Running multiple correlation method” based on the method of running cross correlation (Usoskin *et al.* 1998; Mishra & Tiwari 2003). In the said method, we have calculated the multiple correlation coefficients between different sets of SA parameters and CRI by the following formula

$$R = \sqrt{[(r_{yx_1})^2 + (r_{yx_2})^2 - 2r_{yx_1}r_{yx_2}r_{x_1x_2}]/[1 - (r_{x_1x_2})^2]},$$

where r_{yx_1} is the cross correlation coefficient between dependent variable y and independent variable x_1 , r_{yx_2} is the correlation coefficient between y and independent variable x_2 , and $r_{x_1x_2}$ is the correlation coefficient between independent variables x_1 and x_2 . Here, the different SA parameters have been assumed as independent variables and CRI as dependent variable. Firstly, r_{yx_1} , r_{yx_2} and $r_{x_1x_2}$ have been calculated by the running cross correlation method for every month and then the multiple correlation coefficients have been calculated by putting these values into the formula. As such, the value of R has been obtained for each month of 1955–2004.

3. Results and discussion

For the present study, we have chosen the CRI data observed at Kiel ($R_c \sim 3$ Gv), a middle latitude neutron monitor station. The overall anti-correlation between CRI and different SA parameters is well known, however the magnitude (strength) of correlation varies with time (Gupta *et al.* 2006). To observe these variations and to see the effect of multiple parameters in the modulation of CRI, the running multiple correlation function between CRI and different sets of independent SA parameters (CRI-SSN-SF, CRI-SSN-GSF, CRI-SSN-TA, CRI-SF-CI & CRI-CI-GSF) is illustrated in Fig. 1.

This type of analysis is useful to explain the momentary behaviour of multiple correlation coefficient with respect to time. Unlike the cross correlation coefficient r , which gives an estimate of both the strength and direction of the association, multiple correlation coefficient R , tells only the strength of the association. It is observed from Fig. 1 that the strength of multiple correlation is different for the different phases of particular solar cycle and it varies with time. It is seen that the general behaviour of multiple correlation coefficient for different sets of SA parameters is similar except for cycle 23, which confirms the peculiar behaviour of cycle 23 in relation to previous cycles. Moreover solar cycle 21 also looks peculiar in terms of correlation coefficient among different sets of solar parameters and CRI, which is unexpected and needs further investigation in this direction. It is also evident that the strength of correlation is stronger during ascending and descending phases of solar cycle and it becomes

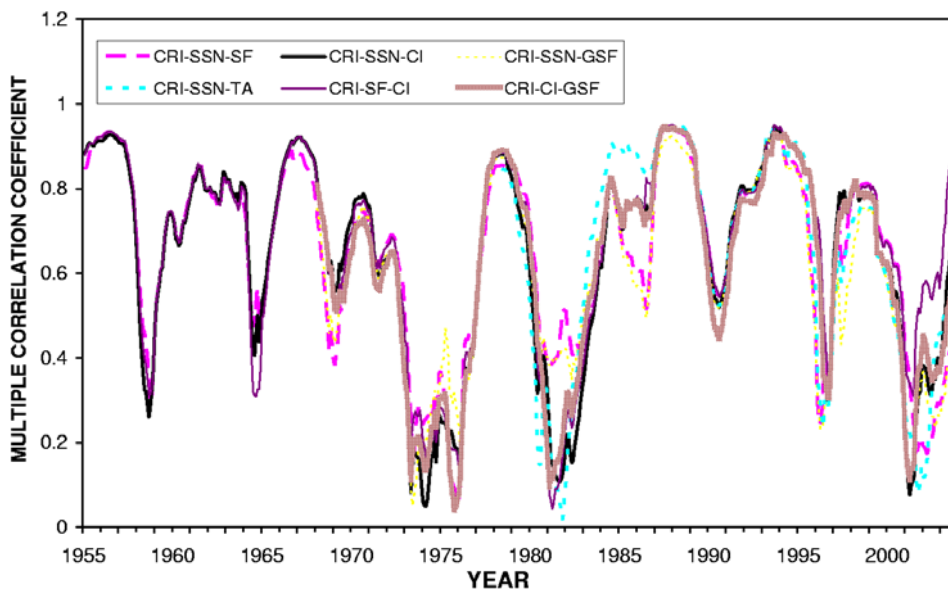


Figure 1. Shows multiple correlation coefficient between CRI and different sets of solar activity parameters during the period 1955–2004.

weaker during maxima and minima of solar cycle, which is in accordance with the results obtained by the running cross correlation method (Gupta *et al.* 2006). Since the variation in sunspot activity which determines the behaviour of a particular cycle depends on solar dynamo process, hence the peculiar behaviour of cycle 23 gives an idea to speculate the behaviour of solar cycle 24. Many authors have speculated the behaviour of solar cycle 24 in contradiction with each other. On the one hand, a group of scientists have speculated it to be exceptionally low (Choudhuri *et al.* 2007) and on the other hand, scientist have urged it to be of high activity (Dikpati *et al.* 2006).

Further, we have compared the results of linear relationship of CRI and SSN (obtained by the running cross correlation method) with the results obtained by the multiple correlation analysis (Fig. 2).

It is obvious from Fig. 2 that there is no significant difference in the results of these two types of analysis throughout the entire period of investigation. The values of multiple correlation coefficients among various sets of SA parameters and CRI falls well within the 95% confidence interval.

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 path of individual particles through the heliosphere is determined by the interplanetary magnetic field lines including drift processes. 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 sign once 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 influence of curvature of interplanetary magnetic field on the transport of cosmic ray should also be considered.

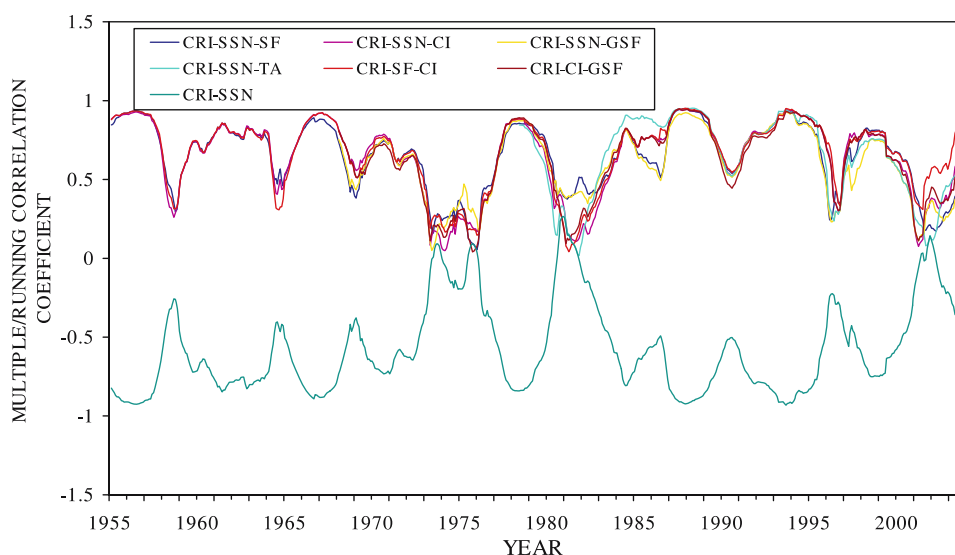


Figure 2. Multiple correlation coefficient between CRI and different sets of solar activity parameters (upper panel) and running cross correlation coefficient between CRI and SSN (lower panel).

4. Conclusions

Based on the observational results presented above, the following conclusions are drawn:

- The strength of multiple correlation is different for the different phases of the solar cycle.
- The strength of multiple correlation is stronger (~ 0.8 – 0.9) for the ascending and descending phases of solar cycle and it becomes weaker (~ 0.2) during the maxima and minima, which is in accordance with the results obtained by the running cross correlation method, where it is found to be ~ 0.8 – 0.9 and ~ 0.1 – 0.2 for both the cases respectively (Gupta *et al.* 2006).
- The overall behaviour of multiple correlation function for different sets of SA parameters is almost similar throughout the period of investigation, except for solar cycle 23 (1996–2005).
- No significant difference in the results of running cross correlation analysis and multiple correlation analysis has been observed, however, some more sets of parameters (interplanetary and heliospheric) such as solar wind velocity, plasma temperature, coronal holes area, etc., are required to generalize these results.
- In the view of even–odd hypothesis, the strange behaviour of cycle 23 (odd cycle) lead us to speculate that solar cycle 24 might be of an exceptional nature as mentioned above.

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Solar and Interplanetary Disturbances causing Moderate Geomagnetic Storms

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Abstract. The effect of solar and interplanetary disturbances on geomagnetospheric conditions leading to 121 moderate geomagnetic storms (MGS) have been investigated using the neutron monitor, solar geophysical and interplanetary data during the period 1978–99. Further, the duration of recovery phase has been observed to be greater than the duration of main phase in most of the cases of MGS. It has further been noted that Ap-index increases on sudden storm commencement (SSC) day than its previous day value and acquires maximum value on the day of maximum solar activity. Generally, the decrease in cosmic ray (CR) intensity and Dst begins few hours earlier than the occurrence of MGS at Earth. Furthermore, negative Bz pointing southward plays a key causal role in the occurrence of MGS and the magnitude and the duration of Bz and Bav also play a significant role in the development of MGS. The solar features H α , X-ray solar flares and active prominences and disappearing filaments (APDFs) which have occurred within lower helio-latitudinal/helio-longitudinal zones produce larger number of MGS. Solar flares seem to be the major cause for producing MGS.

Key words. Solar flares—active prominences and disappearing filaments—sunspot numbers—geomagnetic storms.

1. Introduction

Geomagnetic disturbances are generally represented by geomagnetic storms, sudden ionospheric disturbances (SIDs) and ground level enhancements (GLEs). During the geomagnetic storms, it has been observed that horizontal component of Earth's magnetic field (H) may attain a maximum value in few to several minutes, the rise is quicker at noon ($\simeq 2$ min) than that at mid night ($\simeq 5$ min). The change of the magnetic field during Storm Sudden Commencements (SSCs) (Burlaga and Ogilvie 1969) is step like and oscillatory.

The impact of an interplanetary irregularity causing a sudden change in the geomagnetic components are called SSCs (Smith *et al.* 1986). It is accepted that the impact of the interplanetary irregularity on the magnetopause causes hydromagnetic waves which travel through the magnetosphere towards the Earth as also into the tail and their isotropic modes give the SSC. Almost all the observable features of the solar

atmosphere owe their existence to the sunspot magnetic field (Kane 1976). Hewish & Bravo (1986) found that geomagnetic storms are more associated with coronal holes than the solar flares. Burlaga *et al.* (1987) have concluded that these are associated with either a compound stream or a magnetic cloud which is also confirmed by Wilson (1987). Further, co-rotating interaction regions and some other phenomena has also been observed to cause GMSs.

Recently, it has been shown that CMEs without considering solar flares are the key causal link with solar activity and produce geomagnetic storms (Webb 1995). Data available from Skylab mission suggest that the coronal holes, CMEs, Eruptive Prominences and Disappearing Filaments (EPDFs) have causal link with solar activity and energy emitting regions and they produce geomagnetic storms. In this paper, a detailed analysis of moderate geomagnetic storms (MGS) has been presented and an attempt has been made to understand the relationship between solar and interplanetary causes with MGS.

2. Data analysis

In order to examine the relationship between solar and interplanetary causes with MGS in a statistical way, we have considered 121 MGS during the period January 1978 to December 1999. Because of limited space, only four typical cases of MGS observed have been listed in the Table 1. In the present study, MGS are considered with horizontal geomagnetic field component, $H \leq 250\gamma$ (Pandey and Kohli 1982) and $A_p \geq 20$. The data from neutron monitoring stations have also been analysed. The hourly solar wind velocity (V) data is taken from interplanetary medium data books, available up to period December 1993. These data are measured through IMP-8 satellite (Couzens & King 1986, 1989, 1994). The positions of associated $H\alpha$ solar flares, X-ray solar flares and APDFs have been noted 1 to 6 days prior to the occurrence of MGS at Earth depending upon Solar Wind velocity (V) (Solar Geophys. Data Reps., 1978–99).

3. Salient features of MGS observed

The initial phase of storm starts after abrupt change in B_z component either from N to S or from S to N. The main phase of storm starts after sudden increase in V and the average IMF (B_{av}). There is also abrupt change in B_z component from northward to southward. Not only negative B_z or B_{av} is responsible for development of MGS but also its magnitude and duration contribute significantly (Kane 1977). It is deduced from the plots of all the events that the durations of the initial, main and recovery phases are lying between 20–52 hours, 5–20 hours and 21–30 hours respectively. Further, it is observed in case of individual events that the duration of the recovery phase is always larger than the duration of main phase as is clear from the plot of January 1990.

A peculiar result has been observed during the years 1983 and 1994 when Sunspot numbers (SSNs) decrease very rapidly while number of MGS increases (Fig. 1). It is observed that 48% MGS are produced by eastern X-ray solar flares; whereas 52% MGS are produced by western X-ray solar flares. Further, 78% MGS are associated with X-ray solar flare at the heliographic longitude in the range $0-60^\circ\text{E}$ and $0-60^\circ\text{W}$. Remaining 22% are distributed over the range $60-90^\circ\text{E}$ and $60-90^\circ\text{W}$. Thus, X-ray

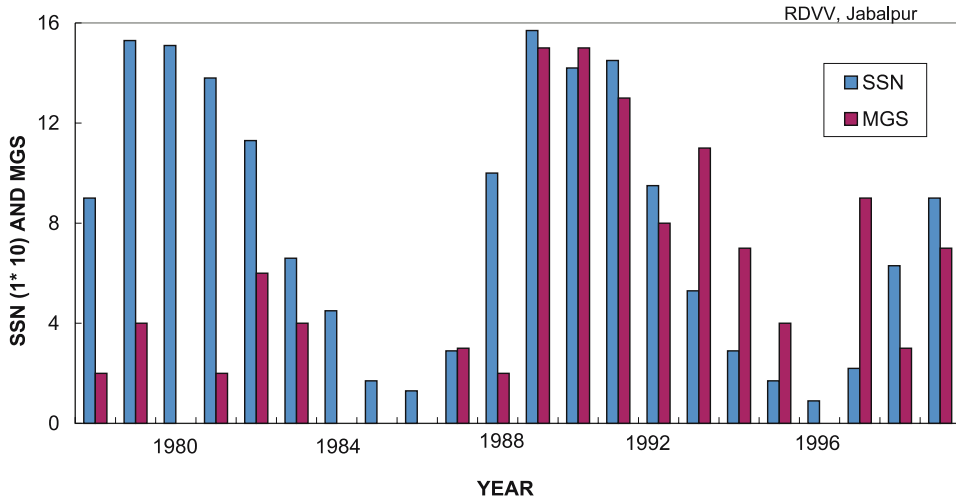


Figure 1. Frequency occurrence histogram of yearly distribution of MGS and their association with SSN.

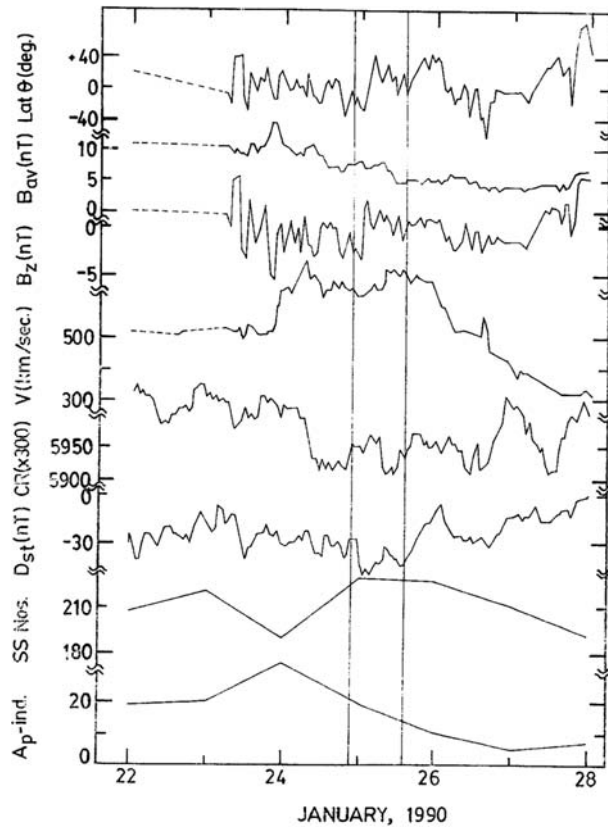


Figure 2. IMF and SWP parameters, i.e., latitude, θ (deg); B_{av} (nT); B_z (nT); V (km/sec); CR intensity; D_{st} (nT); SS nos. and A_p index have been plotted during the period 22–28 January 1990 for the 24th January 1990 event.

solar flare occurring within lower heliographic latitude and longitude are able to produce a strong configuration of closed magnetic field regions which cause fast IP shocks in the interplanetary medium and MGS on the Earth.

On the higher latitude and longitude $H\alpha$, X-ray solar flares do not produce fast IP shocks. There is equal distribution of APDFs in both northern and southern hemispheres; however, in helio-latitude range $0-30^\circ\text{N}$ and $0-30^\circ\text{S}$, 89% of APDFs are associated with MGS. 03 MGS out of 121 investigated, are not associated with any solar features. This shows that some solar features occurring on the back side of the solar disc are also contributing for this cause. Interplanetary Magnetic Field (IMF) and Solar Wind Plasma (SWP) parameters along with CR intensity recorded at Deep River neutron monitoring station, Dst, SSNs and Ap indices have been depicted in Fig. 2 for the MGS of January 1990. Figure 2 also shows the beginning and ending boundaries of interplanetary magnetic cloud affecting these parameters.

4. Conclusions

- In solar cycle 21, 69% of MGS occurred during maximum phase and 31% during minimum phase whereas, in solar cycle 22, 59% of MGS occurred during maximum phase and 41% during minimum phase. Thus, maximum number of MGS have occurred during the maximum phase of solar cycle.
- $H\alpha$, X-ray solar flares, occurred within lower heliographic latitude $0-30^\circ\text{N}$ and $0-30^\circ\text{S}$ and longitude $0-60^\circ\text{E}$ and $0-60^\circ\text{W}$, produce maximum number of MGS. On the higher helio-latitude beyond 40° and helio-longitude beyond 60° ; $H\alpha$, X-ray solar flares do not produce fast IP shocks. A few number of less intense geomagnetic storms are associated with these zones.
- Fast and slow IP shocks are associated with different properties of $H\alpha$, X-ray solar flares and APDFs, e.g., NOAA region, location (helio-latitude/helio-longitude), duration and shape of solar event.
- 54.4%, 30% and 47% MGS are associated with $H\alpha$, X-ray solar flares and APDFs respectively.

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Effect of Magnetic Activity on Ionospheric Time Delay at Low Latitude

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Abstract. The purpose of this work is to investigate the effect of magnetic activity on ionospheric time delay at low latitude Station Bhopal (geom. lat. 23.2°N, geom. long. 77.6°E) using dual frequency (1575.42 and 1227.60 MHz) GPS measurements. Data from GSV4004A GPS Ionospheric Scintillation and TEC monitor (GISTM) have been chosen to study these effects. This paper presents the results of ionospheric time delay during quiet and disturbed days for the year 2005. Results show that maximum delay is observed during quiet days in equinoxial month while the delays of disturbed period are observed during the months of winter. We also study the ionospheric time delay during magnetic storm conditions for the same period. Results do not show any clear relationship either with the magnitude of the geomagnetic storm or with the main phase onset (MPO) of the storm. But most of the maximum ionospheric time delay variations are observed before the main phase onset (MPO) or sudden storm commencement (SSC) as compared to storm days.

Key words. Ionospheric time delay—scintillation—TEC—geomagnetic storm.

1. Introduction

The ionospheric refraction remains a major error source in Global Positioning System (GPS) for real-time applications. After integrating the phase and group refractive indices along the path of the GPS signal, the range obtained between the satellite and the receiver is different from the true geometric range by the amount called ionospheric error. This error is negative for carrier phase pseudoranges and positive for the code pseudoranges (Komjathy 1997). The ionospheric delay is proportional to the number of electron content and inversely proportional to the frequency squared. The total delay contains the dispersive ionosphere, transmitter and receiver delays plus non-dispersive delays due to clocks and the troposphere. Therefore, the change of pseudorange measurement caused by the ionospheric refraction may be restricted to the determination of the Total Electron Content (TEC). The ionosphere causes GPS signal delays due to the TEC along the path from the GPS satellite to receiver.

At equatorial and low latitudes, TEC is subjected to dynamical changes owing to the changes in the electric field. For high precision GPS positioning, the ionosphere effect must be estimated so that the more precise position result could be measured.

Owing to the continuous operation and the large number of world wide receivers, GPS is a powerful tool to investigate ionospheric structures, mainly during magnetically disturbed periods when dynamics and energy dissipation processes in the magnetosphere–thermosphere–ionosphere system become extremely complex (Prolss *et al.* 1995; Fuller-Rowell *et al.* 1997). The strong electric fields that are generated cause significant changes in the ionospheric morphology, producing large delays of the propagation of the GPS signals and an advance in the phase of the carrier. The intensity of a magnetic storm is commonly defined by the minimum Dst value during the magnetic storm, which is equivalent to the maximum Dst magnitude at the main phase. As compared to quiet geomagnetic conditions the daytime ionization enhanced at high latitudes and after that it moved to low latitudes. The ionospheric response to magnetic storms is latitude dependent and also varies with the intensity of the storm and its time of occurrence. The present paper stresses especially on the response of the ionospheric time delay in equatorial regions of the ionosphere as it is directly proportional to the total electron content. The data were used for analysis of ionospheric time delay behaviour during the storm. The objective of this paper is to study the effect of magnetic activity on ionospheric time delay at low latitude regions.

2. Experimental data and methodology

GPS data obtained using code a measurement is only used from January 2005 to December 2005 at Bhopal (lat. 23.2°N, long. 77.6°E). From the processed data, elevation angle and TEC are used to estimate the time delay values at elevation cut off 40°. Here the ionospheric delay used is: $\Delta \text{ion} = 40.3/\text{cf}^2 \text{ TEC}$. In order to clearly identify the ionospheric time delay, we compute the delay for five international quiet (*Q*) and disturbed (*D*) days and during storm conditions.

3. Result

The annual variation of occurrence of ionospheric time delay on quiet and disturbed days during the same period is shown in Fig. 1. It is observed from the figure that the maximum ionospheric time delay occurred in the month of October during quiet period while the maximum delays of disturbed period are observed during the months of winter. The occurrence of ionospheric time delay on *Q*- and *D*-days for the three seasons; winter, summer and equinox months and their annual average are shown in Fig. 2. It is obvious from this figure that, under the geomagnetic disturbances, the maximum ionospheric time delay occurred in equinox while minimum in winter and moderate in summer. Result shows that seasonal dependency is one of the major factors of ionospheric time delay under the effect of magnetic activity during low solar activity.

3.1 Case studies

The analysis represents the ionospheric time delay for all the storms in relation with Dst respectively for one day before and after the Sudden Storm Commencement (SSC)

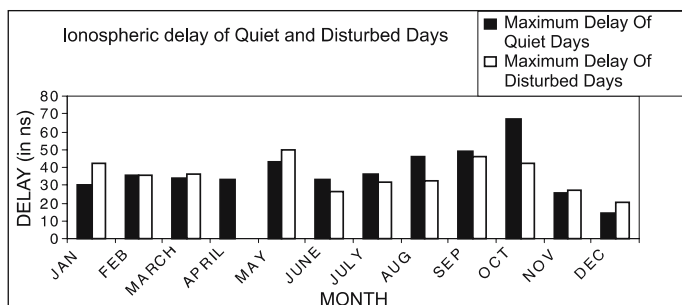


Figure 1. Ionospheric time delay for quiet and disturbed days of 2005.

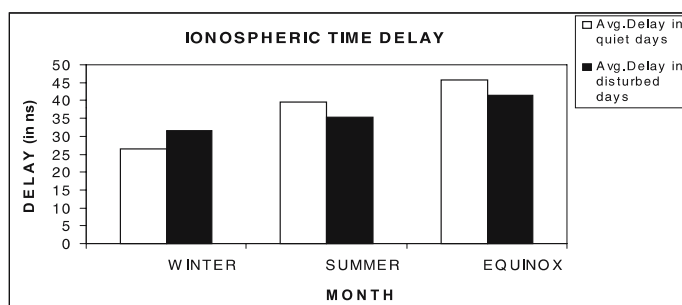


Figure 2. Ionospheric time delay for quiet and disturbed days for the three seasons of 2005.

that describes the ionospheric conditions of Bhopal in response to geomagnetically disturbed conditions for the year 2005.

Case 1. The storm of 7 January 2005, SSC occurred at 0922 UT as mentioned in Fig. 3. The main phase of the storm started about 1500 UT and Dst had its lowest value equal to -96nT on 8 January at 0300 UT. The occurrence of ionospheric time delay was larger on the day before the SSC, i.e., on 6 January 2005. The maximum ionospheric time delay of 31.5173 ns was observed at 0853 UT.

Case 2. On this particular storm day SSC occurred at 1041 UT as mentioned in Fig. 4. The Dst value had its lowest value equal to -96nT at 0300 UT started about 0300 UT on 8 January. The occurrence of ionospheric time delay was larger on the storm day other than the previous day, i.e., on 9 January 2005. The maximum ionospheric time delay was 28.6425 ns at 0925 UT.

4. Summary and discussion

Ionospheric total electron content produces signal delay and refraction that causes range errors that apply to surveillance and GPS navigation. After investigation the result shows that the ionospheric time delay is a function of both season and earth's magnetic field. The maximum ionospheric time delay occurred in equinox during these magnetic days, more significantly at low latitude which is probably due to the

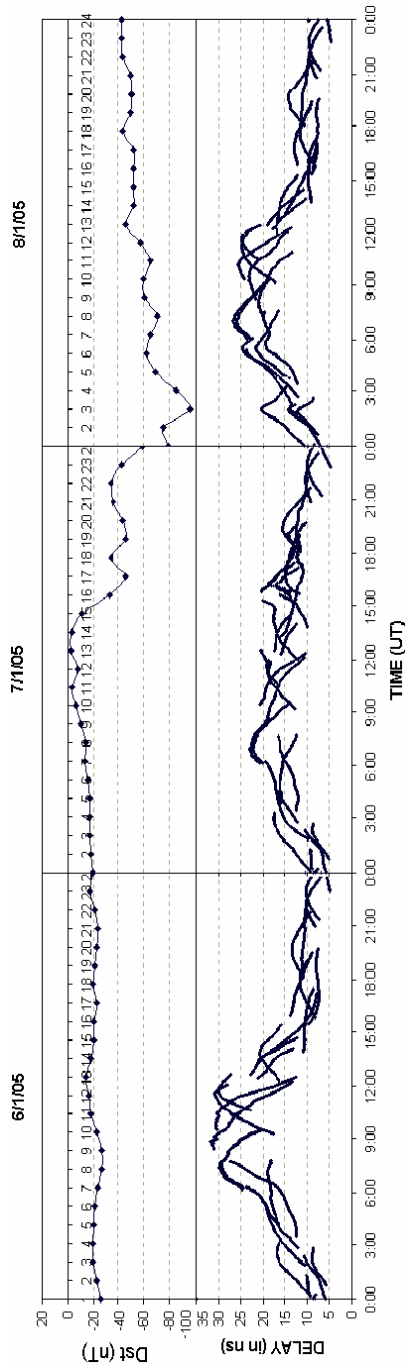


Figure 3. Storm of 7 January 2005 correlated with delay.

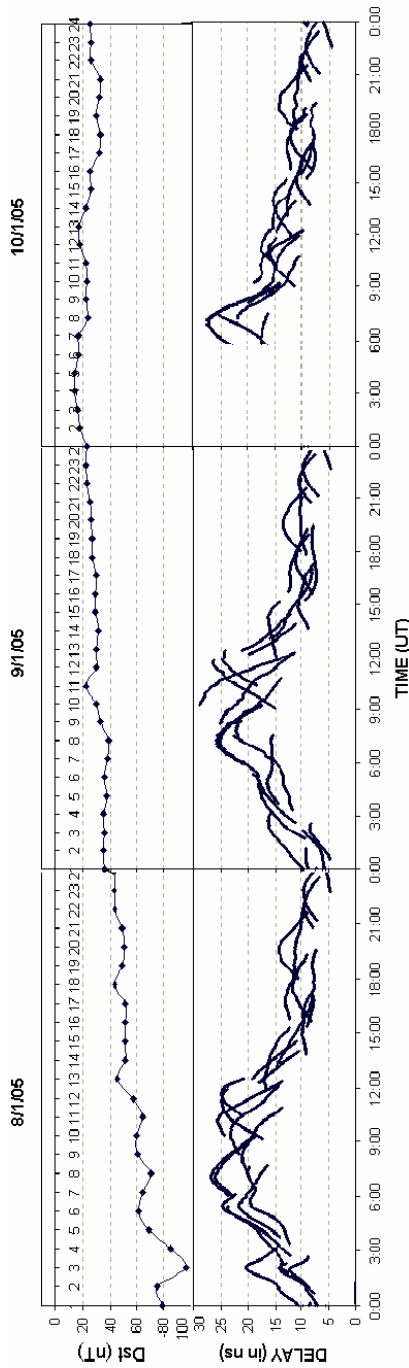


Figure 4. Storm of 9 January 2005 correlated with delay.

downwelling of thermospheric gas. The downwelling of thermospheric gas, mainly caused by storm-induced thermospheric winds (Danilov & Lastovicka 2001), may cause the increase of ionization at low latitudes without any significant change in atom to molecule ratio (Fuller-Rowell *et al.* 1994).

The most favoured mechanisms for the maximum ionospheric time delay are: (a) downward flow from plasmasphere to ionosphere, (b) compression of the plasmasphere by an electric field, (c) upward movement of the F-layer to regions of low loss rate caused either by an electric field or neutral wind, and (d) corpuscular ionization and interhemispheric plasma flow. The relative contribution of different mechanisms could be different at different latitude sectors such as: electric field associated with large pre-reversal height rises during high sunspot equinox at low latitudes and conjugate and protonospheric sources register an increase at high latitudes. During the daytime, equinoctial values, in case of ionospheric time delay is higher than those observed in the winter and summer solstices indicating a semi-annual variation in the whole year of observation irrespective of solar activity condition.

We also observed the maximum ionospheric time delay during magnetic storm condition and the result shows that they do not bear any clear relationship either with the magnitude of the geomagnetic storm or with the main phase onset (MPO) of the storm. But most of the maximum ionospheric time delay variations are observed during quiet days (before MPO or SSC) as compared to storm days. This may be due to changes in electron temperature which have small effect on time delay during disturbed but an appreciable effect during quiet period.

During an ionospheric storm there is heating of the lower part of the thermosphere in the auroral region. The main source of this heating is the joule dissipation of electric currents (Forster & Jakowski 2000; Danilov & Lastovicka 2001) and it should lead to a significant decrease in the atom-to-molecule ratio throughout the entire high latitude atmosphere (Buondanto 1999). Further, the low latitude ionosphere in the Indian region is characterized by large horizontal gradients, intense irregularities and large day-to-day variability. Hence, there is a necessity to thoroughly understand and model the effects of ionospheric time delay on the radio system. And this statistical information would be useful for developing a suitable vertical time delay algorithm for the Indian subcontinent.

5. Conclusion

From the above results we conclude that:

- Under the geomagnetic disturbances, the maximum ionospheric time delay occurred in equinox while minimum in winter and moderate in summer.
- The maximum ionospheric time delay which occurred in the months of equinox, summer and winter was 67.0377 ns, 45.9846 ns and 35.2864 ns during quiet period and 46.0726 ns, 50.0269 ns, 42.6269 ns during disturbed period respectively. While minimum ionospheric time delay was 32.9488 ns, 32.8303 ns, 14.4249 ns during quiet period and 35.9586 ns, 26.2258 ns, 20.1631 ns during disturbed period respectively.
- The average ionospheric time delay in the months of equinox, summer and winter was 45.7875 ns, 39.4716 ns, 26.3558 ns during quiet periods and 41.3250 ns, 35.1827 ns, 31.4715 ns during disturbed periods.

- Result does not show any clear relationship either with the magnitude of the geomagnetic storm or with the main phase onset (MPO) of the storm. But most of the maximum ionospheric time delay variations are observed during quiet days.

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Geomagnetic Field Variation during Winter Storm at Localized Southern and Northern High Latitude

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Abstract. This paper presents the effect of geomagnetic storm on geomagnetic field components at Southern (Maitri) and Northern (Kiruna) Hemispheres. The Indian Antarctic Station Maitri is located at geom. long. 66.03° S; 53.21° E whereas Kiruna is located at geom. long. 67.52° N; 23.38° E. We have studied all the geomagnetic storms that occurred during winter season of the year 2004–2005. We observed that at Southern Hemisphere the variation is large as compared to the Northern Hemisphere. Geomagnetic field components vary when the interplanetary magnetic field is oriented in southward direction. Geomagnetic field components vary in the main phase of the ring current. Due to southward orientation of vertical component of IMF reconnection takes place all across the dayside that transports plasma and magnetic flux which create the geomagnetic field variation.

Key words. Dst—vertical component of interplanetary magnetic field and geomagnetic field components.

1. Introduction

The magnetic field is one of the important properties of the earth. The main magnetic field originates from hydromagnetic processes in the liquid outer core, and its configuration and secular variations also depend upon the structure and dynamic processes in core mental boundary. The geomagnetic field has a regular variation which depends on local time, latitude, season and solar cycle. These variations indicate that the geomagnetic signatures at high latitude are those left by the southern limb of the quiet time *Sq* current loop in the Southern Hemisphere and those left by the northern limb of the quiet time *Sq* current loop in the Northern Hemisphere. The disturbances in the geomagnetic fields are caused by fluctuations in the solar wind impinging on the earth. The most energetic magnetic storms occur during peaks in solar activity and when the vertical component of interplanetary magnetic field (IMF) becomes intimately connected with that of the earth (Gonzales *et al.* 1994). The disturbances may be limited to the high-latitude region unless the vertical component of interplanetary magnetic

Table 1. List of all severe storms that occurred in winter season.

Month	Day	Dst nT	Kp	Ap	IMF	Storm
January 2004	22	−149	7 [−]	64	−19	Severe
February 2004	11	−109	3 ⁺	12	−15	Severe
November 2004	7	−373	8°	50	−51	Severe
	7	−373	8°	50	−51	Severe
	7	−373	8°	50	−51	Severe
	9	−223	9 [−]	119	−39	Severe
	10	−289	9 [−]	161	−28	Severe
January 2005	18	−121	8 [−]	84	−26	Severe
	21	−105	8°	66	−30	Severe

field carried by the solar wind has long periods (several hours or more) of southward orientation with magnitudes (greater than 10–15 nT). In this paper, we examine and compare the winter variations in geomagnetic field components with respect to the variation in global parameter.

2. Data selection

The data used in this study was collected from the Swedish Geomagnetic Observatory “Kiruna” for Northern Hemisphere. For Southern Hemisphere we used Indian Antarctic Research Station “Maitri”. Digital data of magnetic field is collected by fluxgate magnetometer. The vertical component of interplanetary magnetic field discontinuities of the same time (winter season of year 2004–2005) is also detected from ACE satellite. Dst index, Kp and Ap index are taken from World Data Center, Kyoto, Japan. With global indices we compare the local phenomenon to study the variation of geomagnetic field component at localized Northern and Southern Hemispheres.

3. Results

To study the effect of storms on localized stations we have analyzed the winter storms of year 2004–2005. Table 1 shows the list of the most severe storms. Depending upon the Dst value we have categorized the geomagnetic storm in minor ($\text{Dst} > -50$ nT), moderate ($\text{Dst} \geq -100$ nT) and severe ($\text{Dst} < -100$ nT) and we have selected those events in which Dst is < -100 . Before considering the various component of earth’s magnetic field we have considered the vertical component of IMF and compared geomagnetic field component with its orientation at both hemispheres.

Case 1. January 22, 2004. Dst index Kp and Ap sum is shown in Table 1. The initial phase of the ring current started from 1:36 UT to 11:00 UT. The vertical component of IMF oriented to southward direction with value -19 nT for two hours and after that it remains southward for more than 24 hours. At Southern Hemisphere the X component shows the increase of 600 nT. Y component shows the decrease of 600 nT and the Z component shows the decrease of 900 nT. At Northern Hemisphere the X component

decreases upto 600 nT, Y component fluctuates and Z component decreases upto 400 nT. In the main phase of the ring current and southward orientation of interplanetary magnetic field the geomagnetic field components show the maximum variations. Due to southward orientation of vertical component of IMF there is the increase in the central plasma sheet ions. Because of increase in central plasma ions the geomagnetic field varies.

Case 2. February 11, 2004. As its maximum negative excursion value is -109 nT. But the geomagnetic field component does not show any significant variation at Maitri, the X component increases with 300 nT, Y component fluctuates and Z component decreases with values 400 nT. At Kiruna X component decrease of 200 nT and the decrease in Y component is of 200 nT and there is the decrease of 400 nT in Z component.

Case 3. November 7, 2004. There are three storms on 7 November 2004. The maximum negative excursion of the ring current is -373 nT at 06:00 UT. When the vertical component of interplanetary magnetic field turns to southward direction for 12 hours with value of -51 nT at 22:00 UT. At Southern Hemisphere, the X decreases upto 2000 nT, Y component fluctuates and Z component increases upto 1400 nT. At Northern Hemisphere, the X component decreases upto 1700 nT, Y component increases upto 1200 nT and Z component fluctuates.

Case 4. November 9, 2004. The maximum negative excursion of the ring current is -223 nT as shown in Fig. 1. When the vertical component of interplanetary magnetic field (B_z) turns southward direction for 15 hours on 10 November 2004, with value of -39 nT at 21:00 UT, at Southern Hemisphere, the X decreases upto 1600 nT, Y increases upto 1200 nT and Z increases upto 1200 nT. At Northern Hemisphere, X decreases upto 800 nT, Y decreases upto 900 nT and Z increases upto 1000 nT. At Southern Hemisphere, the amplitude of variation is more as compared to the Northern Hemisphere.

Case 5. November 10, 2004. Maximum negative excursion of Dst, Ap, Kp and vertical component of IMF are shown in Table 1. At Southern Hemisphere, there is the decrease of 1525 nT in X component and Y and Z component increases with value 1232 and 1146 nT respectively. At Northern Hemisphere, the variation in X component is -1000 nT and increase in Y component and Z component fluctuates.

Case 6. January 21, 2005. The maximum negative excursion of the ring current is shown in Table 1. The vertical component of interplanetary magnetic field turns to southward direction at 16:00 UT with value of -30 nT. At Southern Hemisphere, the X component decreases upto 800 nT, Y component increases at 900 nT and Z component decreases upto 750 nT. At Northern Hemisphere, the geomagnetic X component decreases upto 800 nT, Y component increases upto 900 nT and Z component increases upto 700 nT.

Case 7. January 18, 2005. Dst, Ap and Kp vertical component of IMF are shown in Table 1. At Southern Hemisphere, the X component decreases upto 1000 nT and Y and Z components increase with values of 300 nT and 400 nT respectively. At Northern Hemisphere, X and Y components decrease upto 600 nT and 200 nT and Z component increases with value 200 nT. Geomagnetic field component also shows smooth behaviour during the recovery phase of the event. Due to southward

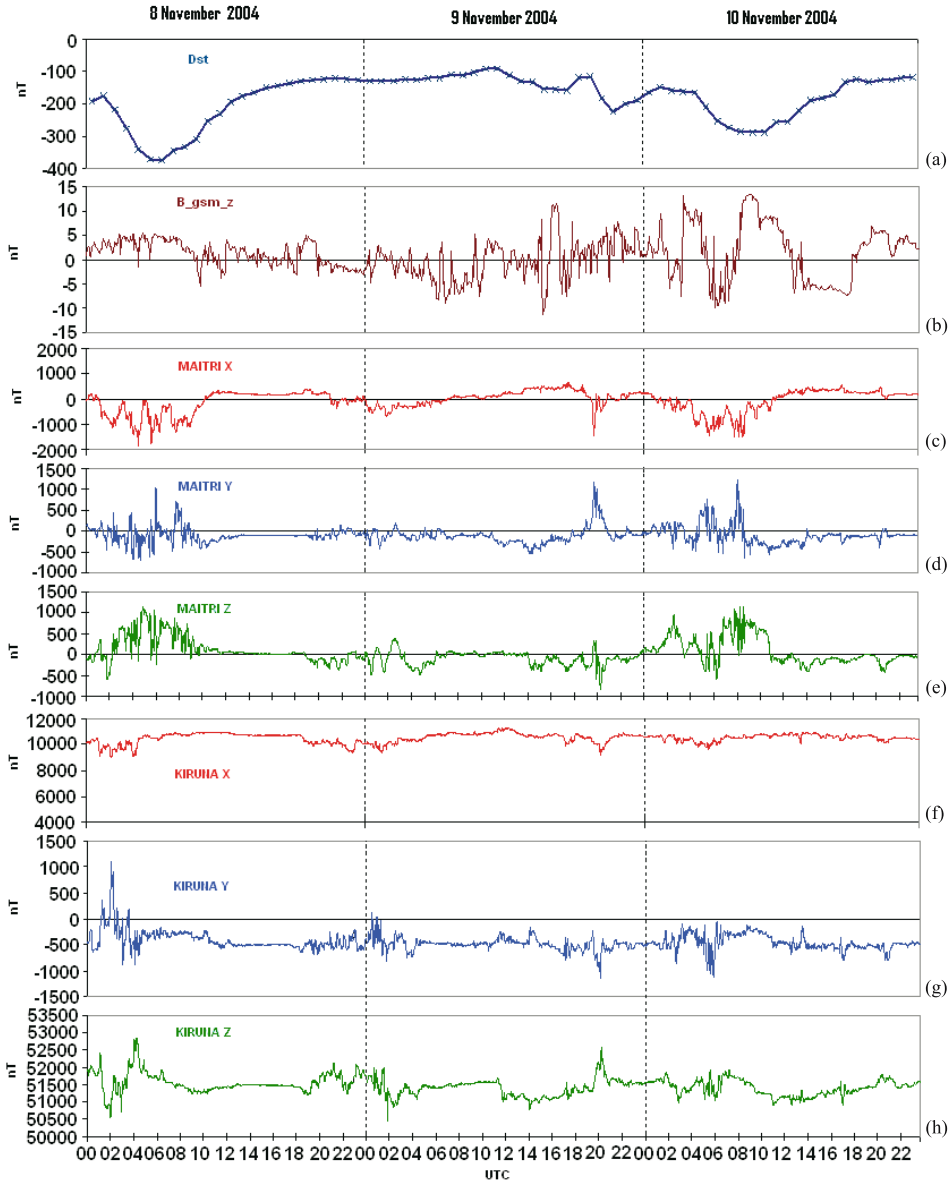


Figure 1. Storm of November 9, 2004, Panel (a) shows the Dst variation, panel (b) shows the vertical component of IMF while panels (c, d, e) show the variation of X, Y and Z components at Maitri and panels (f, g, h) show the variation of X, Y and Z component respectively at Kiruna.

orientation of vertical component, increase in solar wind speed and density, there is an increase in the dynamic pressure which in turn increases the central plasma sheet ions. Because of the increase in central plasma ions the geomagnetic field varies.

4. Summary and discussion

We find a good relation between geomagnetic field components and its causative parameter, i.e., ring current and vertical component of interplanetary magnetic field. When there is a disturbance in the solar wind the current systems existing within the magnetosphere are enhanced and cause magnetic disturbances and storms. The southward orientation of the vertical component of IMF produced substorm activities and large negative Dst variations extending to -350 nT and/or -400 nT, on the ground. On 7 November 2004, the Dst value is observed to be -373 nT. It is observed by (Arnoldy 1971; Foster *et al.* 1971) that the vertical component of (IMF) has been associated with geomagnetic activity. Tsurutani *et al.* (1997) show higher intensities in the ring current and consequently a higher occurrence of geomagnetic storms. The solar wind increases the central plasma sheet ions density. The central plasma sheet ion density controlled mostly by the solar wind proton density and by the northward IMF component. Larger/smaller solar wind ion densities and northward/southward interplanetary magnetic field results in larger/smaller density in the central plasma sheet which in turn results in the hotter/cooler central plasma sheet. According to Ebihara & Ejiri (1998, 2000) the solar wind plasma is supplied to the plasma sheet. Due to the return convection, particles of a boundary transient layer are transported to the inner edge of the plasma sheet and are accelerated. Magnetic field lines of the dayside magnetopause are transported to the tail increasing a magnetic flux. Hence from the above study of all the severe geomagnetic storms of winter season of the year 2004–2005, we observe that

- At Southern Hemisphere the geomagnetic field variation is larger as compared to that of the Northern Hemisphere.
- The geomagnetic field components vary in the main phase of the ring current and when the southward orientation of (IMF) is observed.

In the recovery phase there is smooth variation. In this study we examined winter season storms only, but in future we will analyze the storms during all seasons and their correlation with geomagnetic field components.

5. Conclusion

With disturbed conditions at the front of the magnetosphere, the geometry is unstable and field reconnection occurs with the geomagnetic field and causing acceleration of charged particles in the magnetosphere, where neighbouring but oppositely directed field lines can reconnect, with resulting recoil and acceleration of charged particles in the magnetosphere. A related effect occurs in the tail of the magnetosphere. During the disturbed condition when the vertical component of IMF is southward the maximum convected magnetic energy is integrated over the dayside magnetopause. Due to this, reconnection takes place all across the dayside that transporting plasma and magnetic flux which create the geomagnetic field variation. These variations are more at Southern Hemisphere due to the dayside as compared to the Northern Hemisphere.

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Effect of Interplanetary Magnetic Field and Disturb Storm Time on H Component

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Abstract. A fluxgate digital magnetometer is used to study the variation of magnitude of H component during geomagnetic storm events of April, July and November 2004 at southern subauroral localized region at “MAITRI” (geom. lat. 62°S, long. 52.8°E). We also study the effect of vertical component of interplanetary magnetic field (IMF) on the variation of the magnitude of H component during storm time of April, July and November 2004. Results show that before sudden storm commencement (SSC) time magnitude of H component and IMF show smooth variation but after SSC of first storm of 22 July 2004, the magnitude of the H component shows fluctuations and at 09:00 UT it increases, but during second storm of 24 July 2004, the magnitude of H component indicates large fluctuations and it increases rapidly at 04:00 UT.

Key words. Sudden Storm Commencement (SSC)—magnitude of the H component deviation—Interplanetary Magnetic Field (IMF).

1. Introduction

The magnetosphere responds to the solar wind driving function in a variety of complex ways. Various types of disturbances are produced which can be monitored on the Earth to estimate the level of magnetospheric activity (geomagnetic storm and sub-storm). The association of SSC amplitudes in H with the equatorial electrojet currents has been earlier reported in the literature (Rastogi *et al.* 1964; Rastogi 1978). The geospheric environment is highly affected by the Sun and its features such as, X-ray-solar flares, active prominences and disappearing filaments (APDFs), coronal holes, Coronal Mass Ejection (CME), etc., which are responsible for some large/small geomagnetic storms (Gonzalez *et al.* 1994). This paper discusses three successive geomagnetic storm variations of the horizontal (H) component deviation of the geomagnetic field at subauroral zone with the time of the day, season and sudden storm commencement (SSC). The daily variation of H showed a large peak observed around night time at “Maitri”, Antarctica.

2. Data selection and methodology

A ground based digital fluxgate magnetometer is used to measure the earth's magnetic field at southern subauroral localized region "MAITRI" (geom. lat. 62°S , long. 52.8°E). Global geomagnetic Dst index is taken from the World Data Centre Service. Vertical Component of Interplanetary Magnetic Field values are taken from the Advance Composition Explorer (ACE) Satellite Data Service of World Data Centre. Days that are used for this work are successive storm days along with the preceding and succeeding days of equinox month of April 2004, summer month of July 2004, and winter month of November, 2004 and these storms occur after every alternate day. To study the H component deviation during storm time we have first calculated the average of quiet days as a reference of the storm month. With respect to that we have taken the deviation of H component of storm time and studied the effect of vertical component of the Interplanetary Magnetic Field (IMF) on it.

3. Result

08–13 April 2004. There are three panels (I–II–III) in Fig. 1. First panel shows the Dst of successive magnetic storm while second and third panels show the variation in H component deviation and Interplanetary Magnetic Field respectively. Graphs show the results of three magnetic storms dated 9 April, 10 April and 12 April, 2004. The upper panel of Fig. 1 shows the Dst variation for three magnetic storm. First storm began at 02:33 UT (08:20 LT) on 9 April 2004. The second storm began at 20:10 UT (25:40 LT) on 10 April, 2004 and third at 18:17 UT (23:47 LT) on 12 April 2004. Before SSC magnitude of H component deviation and IMF show smooth variation but after sudden commencement of first storm on 9 April 2004, the magnitude of H deviation shows fluctuations and at 01:00 UT it increases small and at 05:00 UT it sharply increased to 800 nT whereas IMF moves towards south and shows the value of 10 nT.

21–27 July 2004. Figure 2 has three panels (I–II–III). The first panel shows the Dst of magnetic storm while second and third panels, show the variation in H component deviation and IMF (Bz) respectively. Graphs show the result of three successive magnetic storms dated 22 July, 24 July and 26 July, 2004. The upper panel of Fig. 2 shows the Dst variation for three magnetic storms. First storm began at 10:36 UT (16:06 LT) on 22 July 2004, the second storm at 06:15 UT (11:45 LT) on 24 July 2004 and the third storm at 2:49 UT (29:29 LT) on 26 July 2004. Before SSC magnitude of H component deviation and IMF show smooth variation, but after sudden commencement time of first storm of on 22 July 2004, the magnitude of H deviation shows fluctuation and at 09:00 UT it increases little, but for the second storm of 24 July 2004, the magnitude of H deviation indicates fluctuation and at 04:00 UT it increases greatly. The magnitude of H component shows fluctuation and at 13:00 UT it sharply increased upto 1600 nT on the night of 27 July. The 27 July 2004 was a severe magnetic storm. Accompanying this the deviation in magnetic field depends on IMF and magnetic storm.

06–12 November 2004. The event of 7 November storm was one of the severe magnetic storms of the year 2004. Figure 3 panel (I–II–III) shows the three successive magnetic storms of 6–12 November 2004. The first panel shows the Dst of magnetic storm while second and third panels show the variation in H component deviation and IMF respectively. First storm began at 02:56 UT (08:26 LT), then at 10:52 UT (16:32 LT) and finally at 18:28 UT (23:58 LT) on 7 November 2004. The second storm

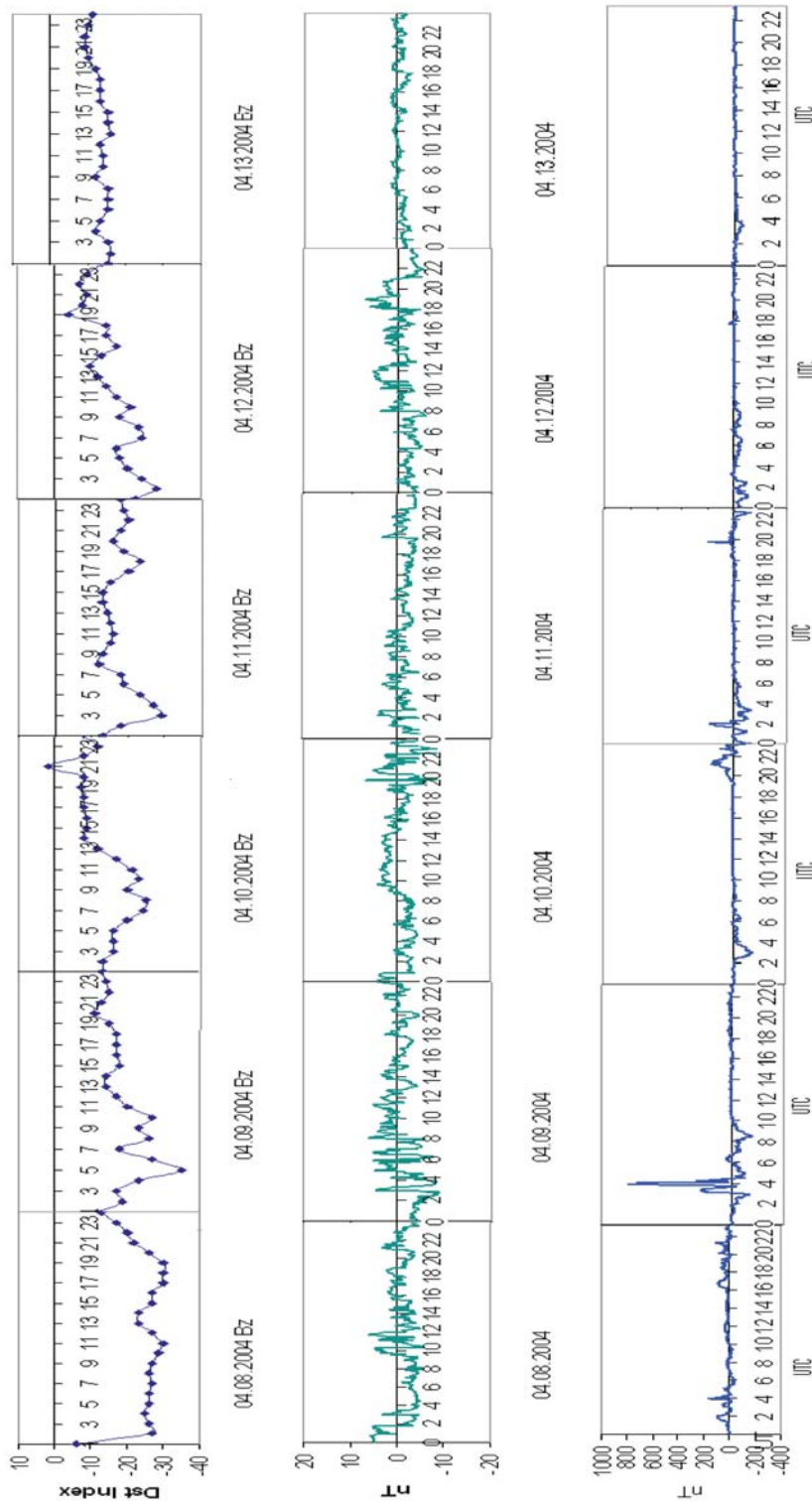


Figure 1. The three panels (from top to bottom) successively show the Dst index on y axis, IMF on y axis, deviation of H component on y axis and time on x axis in UTC which is common to all panels for the storm of 8–13 April 2003.

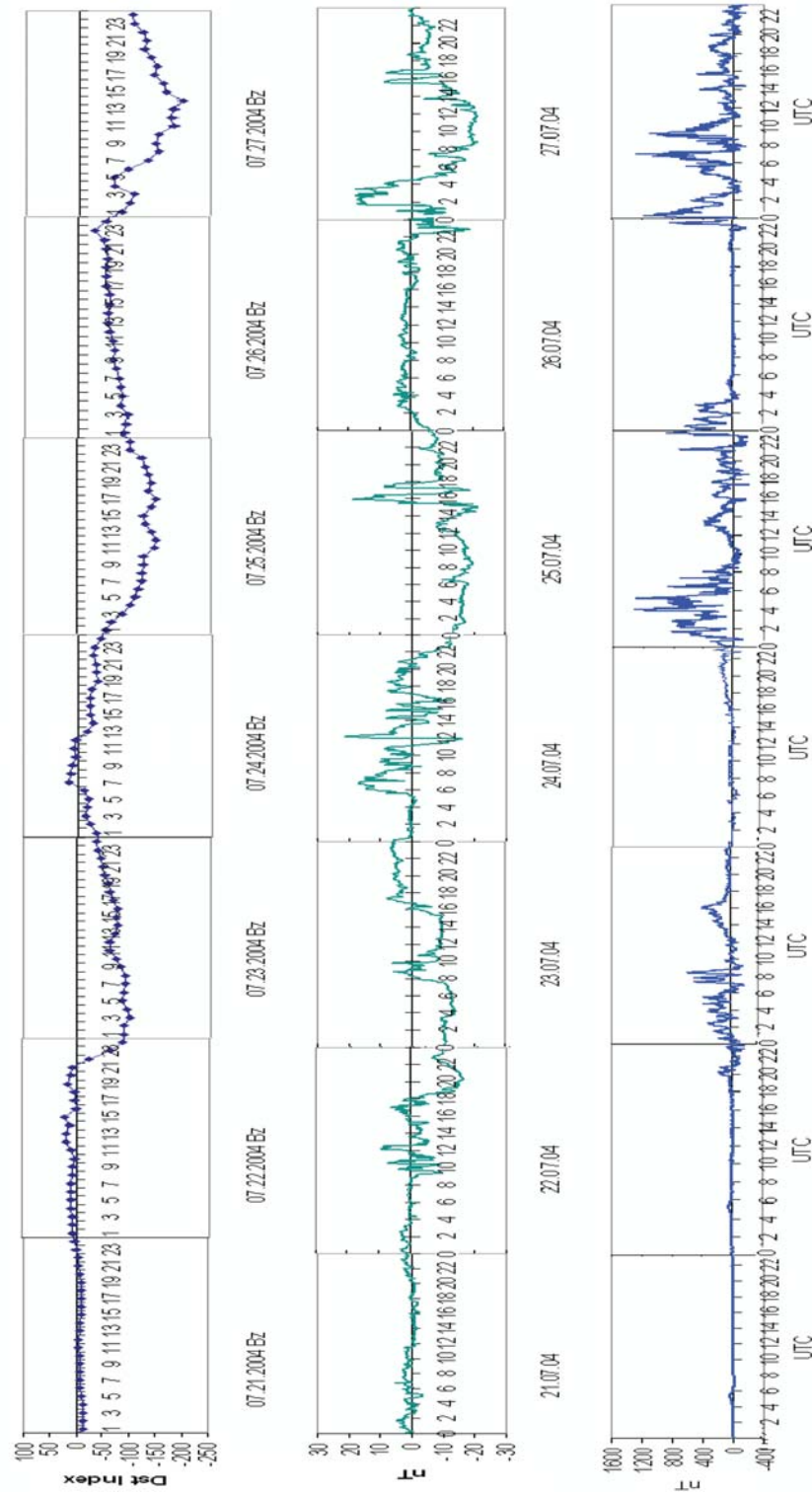


Figure 2. The three panels (from top to bottom) successively show the Dst index on y axis, IMF on y axis, deviation of H component on y axis and time on x axis in UTC which is common to all panels for the storm of 21–27 November 2003.

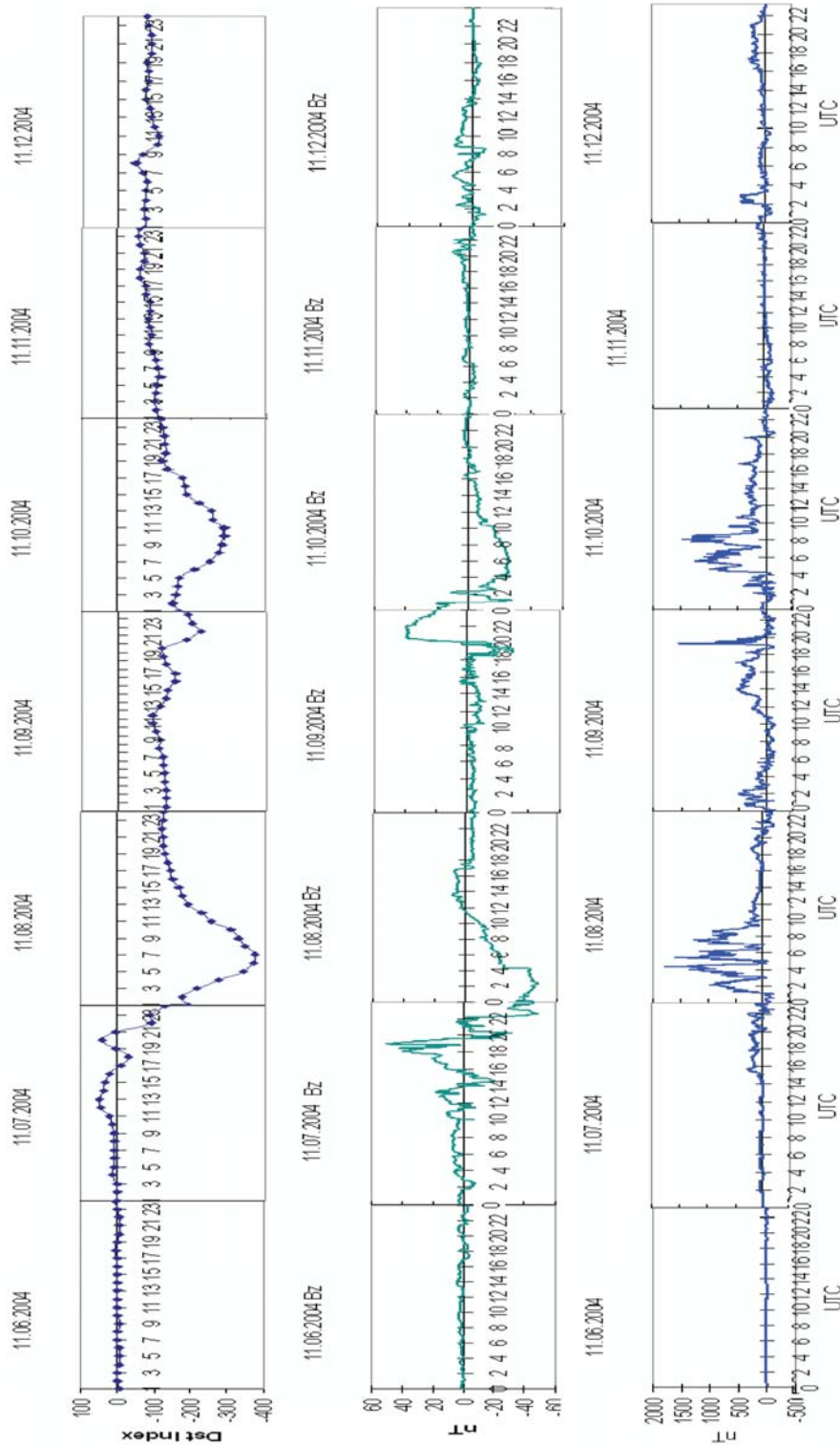


Figure 3. The three panels (from top to bottom) successively show the Dst index on y axis, IMF on y axis, deviation of H component on y axis and time on x axis in UTC which is common to all panels for the storm of 6–12 November 2003.

began at 09:31 UT (15:01 LT) on 9 November and the third storm began at 17:10 UT (22:40 LT) on 11 November 2004. Before SSC magnitude of H component deviation and IMF show smooth variation, but after sudden commencement time the magnitude of H deviation shows fluctuation and at 01:00 UT it sharply increased to 1878 nT on the night of 8 November 2004.

4. Conclusion

The study of southern hemispheric observations of the most energetic solar event of cycle 23 present a very peculiar situation in which there are very strong or the strongest X-ray-solar flares or CME. We observe that even when Bz was positive no significant geomagnetic activity is observed in Earth's magnitude of H component deviation. It is evident that most of the geomagnetic storm activities are associated with negative Bz at subauroral zone (Maitri, Antarctica). Thus either this study had some thing special which needs to be understood or Bz is an essential requirement for generation of strong geomagnetic successive storms and sub-storms at subauroral zone. More studies with other strong solar events are necessary to ascertain these findings. Results show that for the event of 8 November 2004 magnitude of H component shows maximum deviation when Dst shows maximum negative excursion of -373 nT with respect to IMF southward for a prolonged period of time. We observe that oscillations of H deviation in the earth's magnetic field at ground in different seasons is found to be different but the variations is observed alternate days of like waveform with the Sun's IMF after SSC time. The present result shows that the ionospheric conductivity over the magnetic equator and high latitude is directly proportional to the magnitude of H component deviation of the earth's magnetic field. For further studies we can examine the response to impulsive pressure variations in the solar wind.

Acknowledgements

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Geo-effectiveness of CMEs

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Abstract. Coronal Mass Ejections (CMEs) are important phenomena in coronal dynamics causing interplanetary signatures (ICMEs). They eject large amounts of mass and magnetic fields into the heliosphere, causing major geomagnetic storms and interplanetary shocks. Geomagnetic storms are often characterized by abrupt increases in the northward component of the earth's field, called sudden commencements (SSC) followed by large decreases of the magnetic field and slow recovery to normal values. The SSCs are well correlated with IP shocks. Here a case study of 10–15 February 2000 and also the statistical study of CME events observed by IPS array, Rajkot, during the years 2000 to 2003 and Radio Astronomy Center, Ooty are described. The geomagnetic storm index Dst, which is a measure of geo-effectiveness, is shown to be well correlated with normalized scintillation index 'g', derived from Ooty Radio Telescope (ORT) observations.

Key words. Coronal mass ejection—interplanetary medium and normalized scintillation index.

1. Introduction

Coronal Mass Ejections (CMEs) are the most energetic and the largest phenomena associated with the eruption of plasma and magnetic field from the Sun. These are transient expulsions of coronal plasma defying the gravitational field of the Sun. Each CME drains the solar mass in the range of 10^{11} to 10^{13} kg (Webb 1995). CMEs are the major solar drivers of space weather, including non-recurrent geomagnetic storms (Tsurtani *et al.* 1988) and solar energetic particle events (Tsurtani & Lin 1985). CMEs originating from close to the disk center significantly perturb earth's environment and they directly impact the earth (Gopalswamy 2006). Such earth-directed CMEs are the major cause for the severe geomagnetic storms. This paper reports typical geoeffective CMEs observed by IPS array at Rajkot and Ooty and also an important statistical analysis to determine effectiveness of interplanetary parameters for the modeling of geo-magnetic storms and their prediction.

2. Data and results

The interplanetary scintillation observations reported in this study have been obtained from the Rajkot IPS array and Ooty Radio Telescope (ORT) at 103 MHz and 327 MHz respectively. The details of these telescopes and data analysis procedures are given by Vats & Deshpande (2004) and Manoharan (2003) respectively. IPS is a technique for studying the propagation of the solar disturbances from ground-based observations of radio sources (Rickett 1975; Wantanabe & Kakinuma 1984). The scintillation is measured by a statistical index S_4 , which is defined by Briggs & Parkin (1963) as the normalized root-mean-square deviation of signal intensity:

$$S_4 = \frac{\langle (1 - \langle I \rangle)^2 \rangle^{1/2}}{\langle I \rangle} \quad (1)$$

and the normalized scintillation index, g , is derived from long term observations of the interplanetary scintillation by IPS array at Rajkot and Ooty Radio Telescope.

During the years 2000–2003, IPS-Rajkot recorded 25 events, which produced enhancements in scintillation index associated with CME (Jadeja 2007). The interplanetary signatures of one of these, observed during 10–15 February, 2000 are clearly seen in Fig. 1. Here the arrival of the IP shock is characterized by a solar wind velocity increase from ~ 400 km/s to ~ 600 km/s; IMF Bz became negative (-10 nT); ~ 8 fold increase in ram pressure (Pram) and 3-fold increase in density.

For the 10–15 February 2000 event, the IPS sources observed by ORT and IPS array at Rajkot showed enhancements in g -values. The variation of g in ORT and Rajkot data is shown in Fig. 2. Both the temporal variations clearly show the effect of the passage

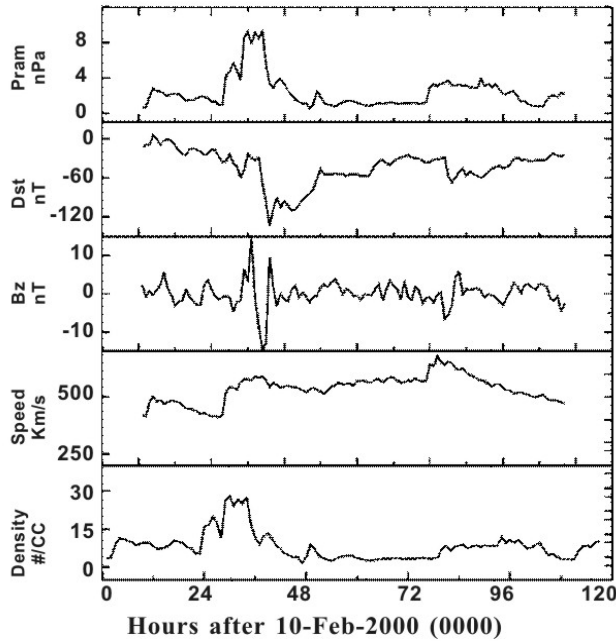


Figure 1. Variation of solar wind and geo-magnetic parameters during 10–15 February 2000.

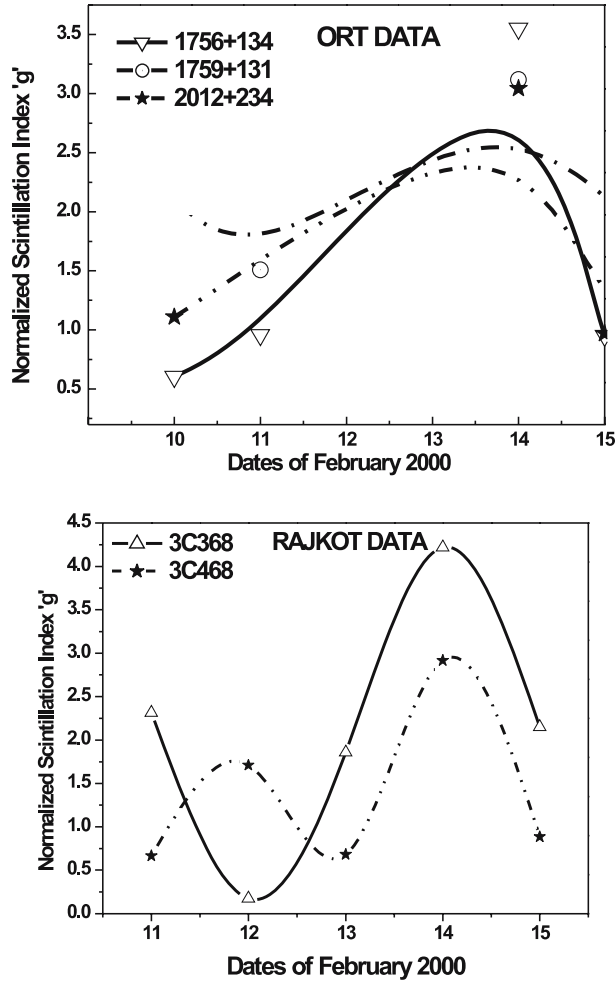


Figure 2. Variation in 'g' observed by ORT during February 10–15, 2000 (top) and by IPS Rajkot during February 11–15, 2000 (bottom).

of the CME plasma cloud or interplanetary shock through the IPM. Three radio sources observed by ORT showed increased g values in the range of 2–2.5 (Fig. 2, top), while two radio sources at Rajkot recorded even higher values of g (2.5 and 4) during this event (Fig. 2, bottom).

Jadeja (2007) reported the statistics of 25 enhanced events in scintillation index observed at Rajkot, of these 2 were associated with severe geomagnetic storms (Dst_{min} between -200 nT and -350 nT); another 2 were associated with strong storms (Dst_{min} between -100 nT and -200 nT); 7 were associated with moderate storms (Dst_{min} between -50 nT and -100 nT) and 10 were associated with weak storms (Dst_{min} between -30 nT and -50 nT). There were 4 events which showed no apparent relation to any storm conditions.

In order to investigate key interplanetary parameters which might be responsible for producing geomagnetic storms, we have selected 30 large CMEs (angular

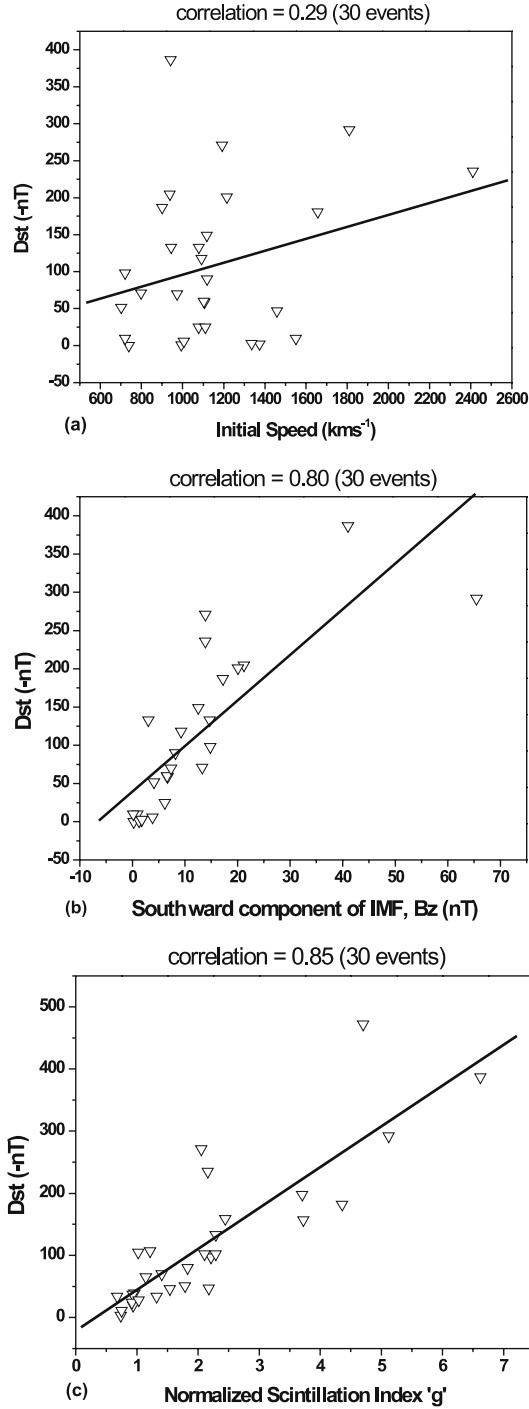


Figure 3. Correlation plots of geo-magnetic index Dst and 3 interplanetary parameters, namely: (a) initial speed of CME (km s^{-1}) measured by LASCO observations, (b) southward component of IMF, B_z (nT) measured by ACE satellite and (c) normalized scintillation index 'g' measured by the IPS observations of radio sources at Ooty.

width $> 150^\circ$) during the period 2000 to 2003. We have used the Dst index to represent the strength of the associated geomagnetic storm (geo-effectiveness) and studied its correlation with 3 IP parameters namely initial CME speed, Bz and g . These are shown in Fig. 3.

Figure 3(a) and (b) respectively show the correlation of Dst with the initial speed of CME as derived from LASCO observations and southward interplanetary magnetic field component Bz measured at 1 AU by ACE satellite. Correlation of Dst with normalized scintillation index ' g ' (which is obtained from Ooty Radio Telescope) is shown in Fig. 3(c). These results indicate that the correlation of Dst with the initial speed of CME, Bz and g are 0.29, 0.80 and 0.85 respectively. From this it is apparent that initial speed of CME has a very poor role in producing geomagnetic storms, whereas g has the highest correlation with Dst, indicating that this parameter is most useful in predicting geomagnetic storms.

3. Discussion and conclusions

Even though Jadeja (2007) found that there are some enhancements in scintillation measurements which do not have a direct relation to the planetary magnetic index, however, the case study shown in Fig. 2 shows quite a large enhancement in g values of both the datasets (ORT and Rajkot). There are few cases of CMEs where moderate to severe deceleration process happens before their arrival at 1 AU, this may result into no apparent effect on the geomagnetic field and its indices. The case study and the statistical results reported here are in good agreement with earlier findings (Lindsay *et al.* 1999; Gopalswamy *et al.* 2000; Manoharan *et al.* 2004) which showed that most of the CMEs having an initial speed greater than ambient solar wind are decelerated. However, few cases of acceleration even when the initial speed is higher than ambient solar wind speed may be due to CME–CME interaction and/or by the energy gained at the time of CME onset and stored within the CME (Manoharan 2006). The IPS measurement in the form of g appears to be a better interplanetary parameter for modeling and prediction of geo-magnetic storms.

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Relationship between Interplanetary (IP) Parameters and Geomagnetic Indices during IP Shock Events of 2005

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Abstract. In the present study, we investigate the possible relationship of IP parameters of solar wind and interplanetary magnetic field with ground-based geomagnetic indices. To carry out the study, we take all the IP shock events listed by Proton Monitor onboard Solar and Heliospheric Observatory (SOHO) during 2005, and plot the time variations of all the IP parameters and geomagnetic parameters (± 5 days), centered at the shock arrival time. Next, we obtain scatter plots of absolute values of solar wind parameters such as V_{sw} , N_{sw} and Interplanetary Magnetic Field (IMF) components B_x , B_y , B_z and total B with the values of geomagnetic parameters such as Dst , K_p indices, dayside Magnetopause (MP) distance and Cosmic-Ray Neutron Monitor count (CRNM). The scatter plots show that before the IP shock, the pattern is random with no clear relationship. Following the shock, a clear pattern emerges with a type of relationship being seen – clear for SHARP shocks and less clear for DIFFUSE shocks. A total of 10 shock events for 2005 have been studied. Typical examples of this behaviour are the shock events of January 21, 2005 and May 15, 2005.

Our study suggests a definite correlation between changes in the solar wind and interplanetary magnetic field parameters and ground-based geomagnetic response. We are trying to obtain quantitative relationships between these for shock events of 2005.

Key words. IP parameters—IP shocks—solar—terrestrial relationship.

1. Introduction

Magnetic field and particle emissions from the Sun are mainly responsible for changes in the electromagnetic and energetic particle environment of earth. An active Sun spews out concentrated particle and field energy into interplanetary space (IP), and manifestations of these have been studied by many researchers (to mention a few, Cargill 2000; Lepping *et al.* 1990; Gopalswamy *et al.* 2004; Bothmer & Schwenn 1998). Propagation of these emissions through IP space and their entry into earth's

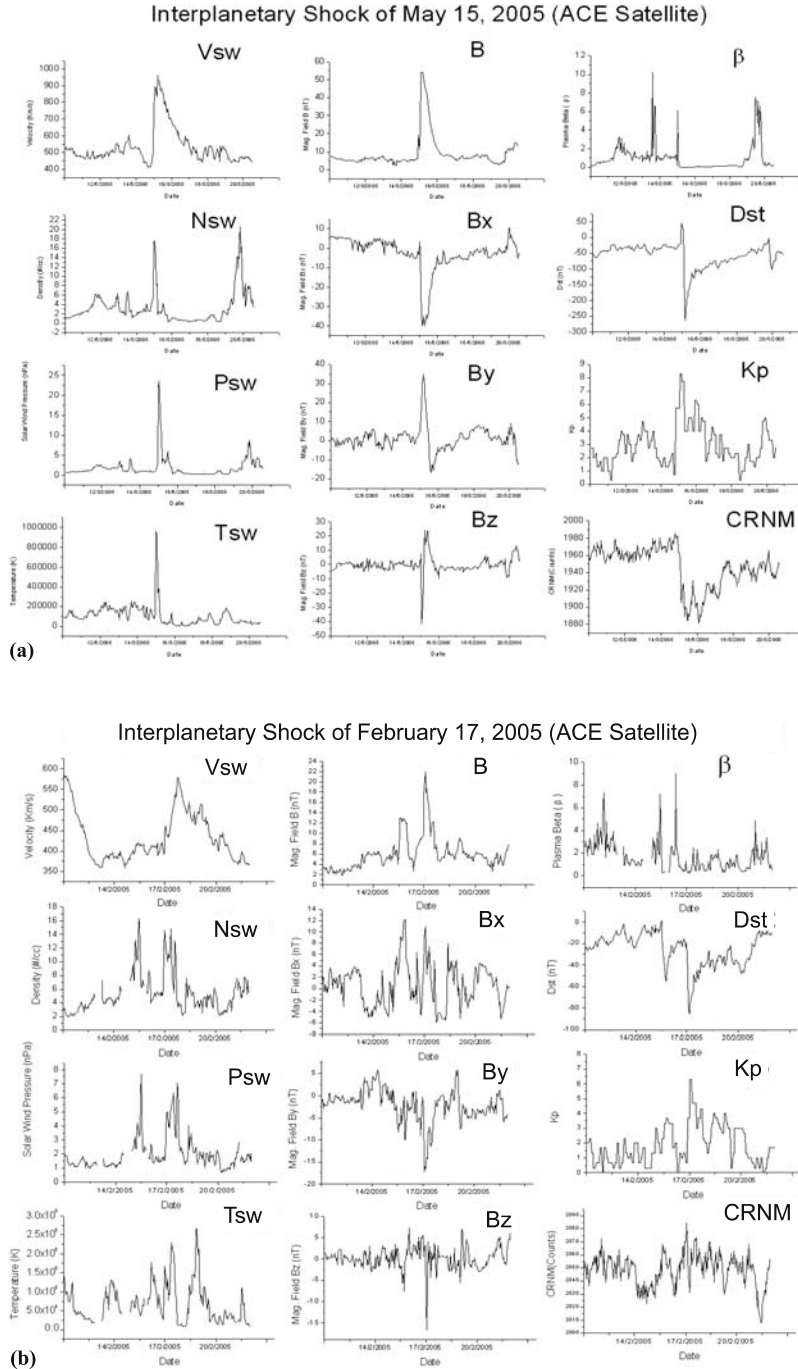
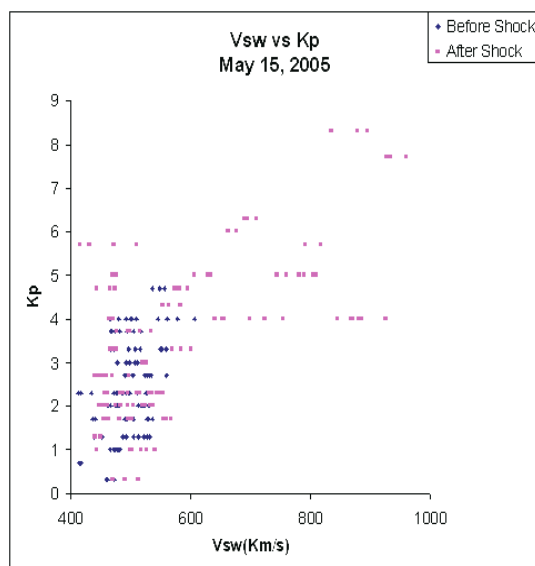


Figure 1. (a) Comparison of solar wind and IMF parameters for ‘sharp’ shock event of 15 May, 2005, and (b) ‘diffuse’ shock event of 17 February 2005. Note the sharp rapid rise and clean drop in these parameters for ‘sharp’ shock as compared to the gradual rise and oscillating variations in the parameters for ‘diffuse’ shock. Note that the β factor and the geomagnetic parameters Dst, Kp and CRNM (Cosmic Ray Neutron Monitor) also reflect the ‘sharp’ and ‘diffuse’ characteristics seen in causative solar wind and IMF parameters.



(a)

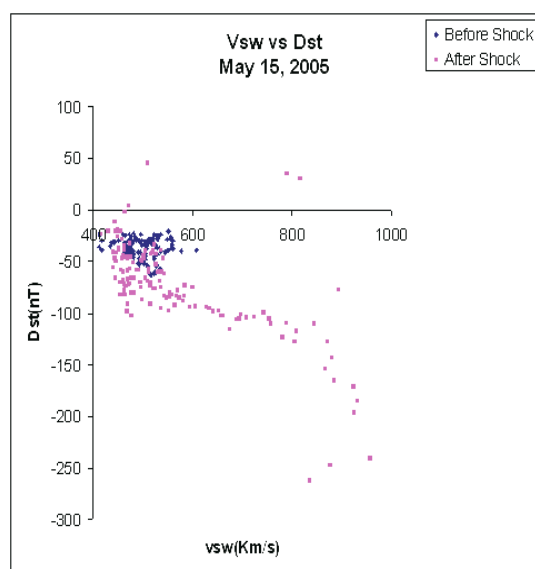
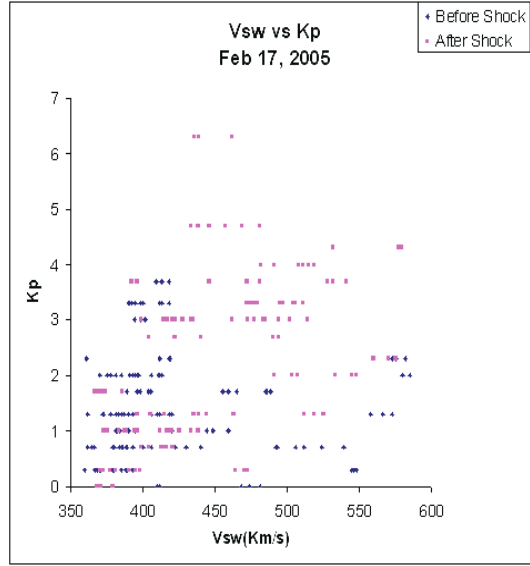


Figure 2(a). Scatter plots of Vsw vs. hourly values of the geomagnetic indices Kp and Dst for ± 5 days around the day of sharp shock (15 May 2005). Note the clustering of points before the shock; after the shock a linear-to-curvilinear relationship is seen for both Kp and Dst.

environment results in various space weather changes, including the phenomenon of the geomagnetic storm and its manifestations (Chapman 1936). Since the 1960s, orbiting and geostationary satellites have made it possible to study the linkage between solar emissions, their manifestations in IP space, and geomagnetic storm phenomena (cf. amongst others Adachi *et al.* 2006; Oh and Yu 2004; Maltsev & Rezhenov 2003; Murayama 1982).



(b)

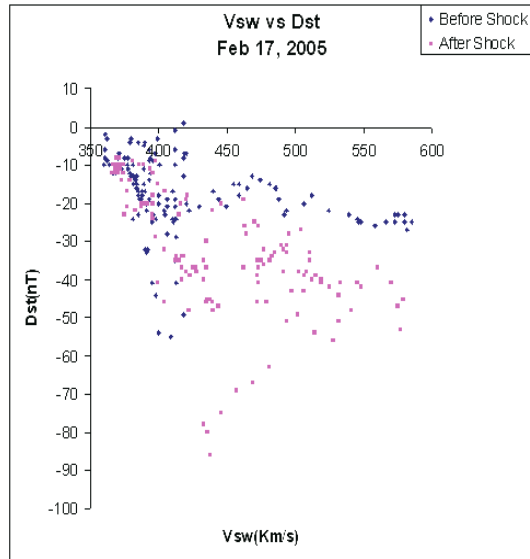
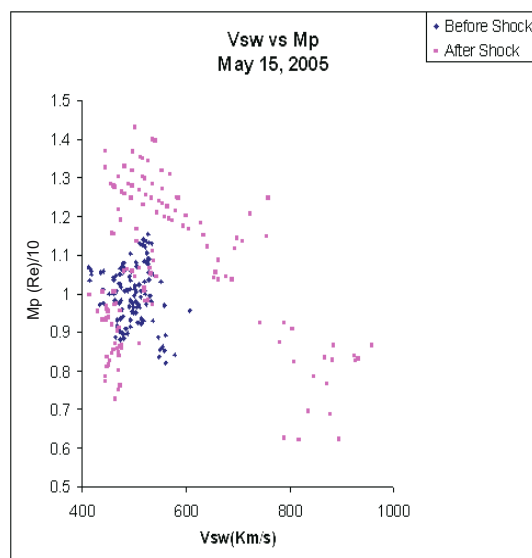


Figure 2(b). Same as for 2(a), but the day is one of a ‘diffuse’ shock (17 February 2005). Note the considerable scatter in relationship between Vsw and Dst and Kp before and after the shock in contrast to ‘sharp’ shock case of Fig. 2(a).

2. Data presentation

For the present work we have studied 19 IP shocks in 2005. In this work, we are studying IP parameters (i.e., solar wind velocity (V_{sw}), solar wind ion density (N_{sw}), solar wind dynamic pressure (P_{dsw}), total interplanetary magnetic field B along with its



(a)

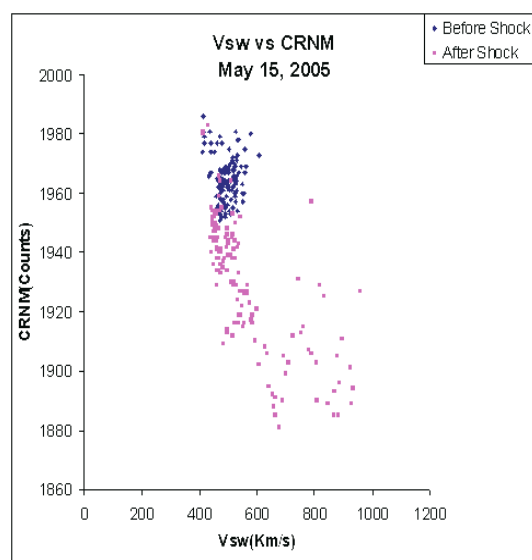
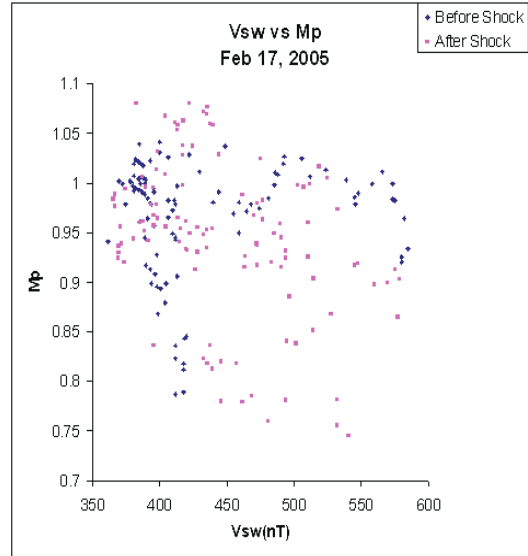


Figure 3(a). Scatter plots of V_{sw} versus hourly values of CRNM and M_p (the Magnetopause distance from centre of the earth) before and after the sharp shock (15 May 2005). Note the clustering of points before the shock. After the shock onset, the M_p moves earthwards as V_{sw} increases. Similarly, CRNM values decrease rapidly with increasing V_{sw} , showing the presence of Forbush-type decrease in cosmic ray protons.

north-south component, B_z , dayside magnetopause distance (MP) and the geomagnetic indices (i.e., Dst , K_p and CRNM count). The data are taken from ACE, OMNI and SOHO satellites. All the data are hourly averaged. The absolute values of all the



(b)

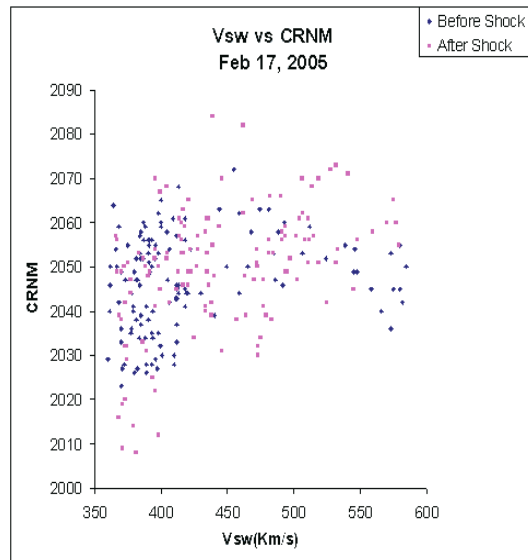
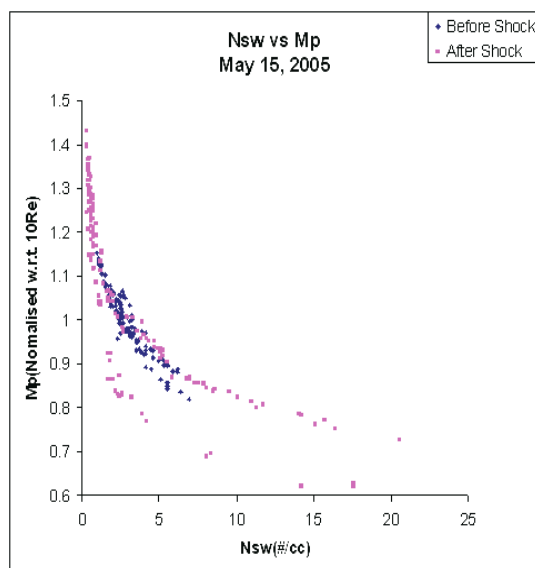


Figure 3(b). Same as for Fig. 3(a), but results are for diffuse shock day 17 February 2005. In this case scatter of Kp values both before and after the shock is considerable for both CRNM and MP showing no clear trend as was seen in Fig. 3(a). In the case of Mp, there is more scatter before the shock onset; after the shock, despite scatter of points, the trend of Fig. 3(a) for MP to decrease with increasing Vsw is seen.

parameters are taken for comparison except for dayside magnetopause distance. In case of magnetopause distance, the values are normalized with respect to $10 R_E$. CRNM counts are taken from Beijing station (www.spidr.ngdc.noaa.gov). The values are taken



(a)

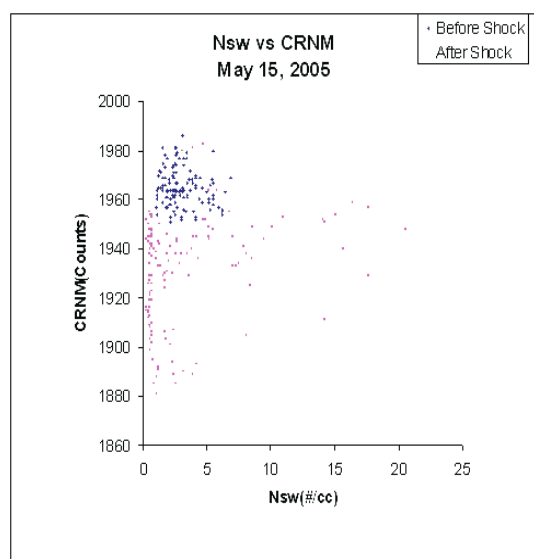
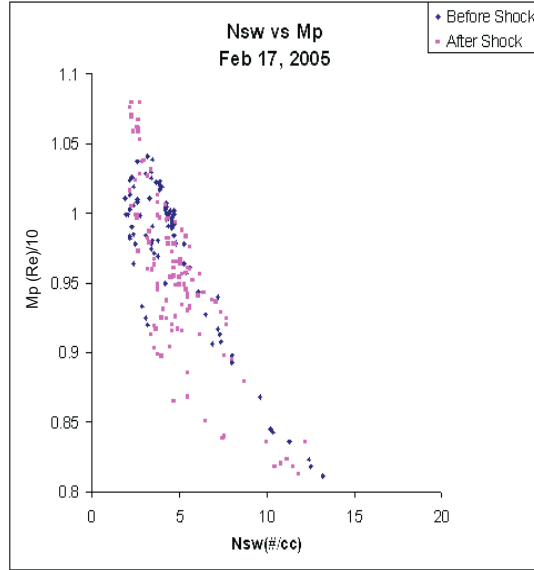


Figure 4(a). Scatter plots of Nsw vs. MP (magnetopause distance from earth) and CRNM (count at Beijing, China) for the sharp shock event of 15 May 2005. Note how the MP moves rapidly after the shock towards earth, first very rapidly with increasing Nsw, then more gradually, i.e., an exponential decrease to values of MP as low as 6Re. The CRNM plot shows that Nsw is not a controlling parameter, as there is no relationship between the two parameters. (MP distance is computed from the balance between the plasma pressure of the solar wind and the magnetic pressure of the geomagnetic field. The y-axis shows the MP distance normalized w.r.t the normal distance 10 Re).



(b)

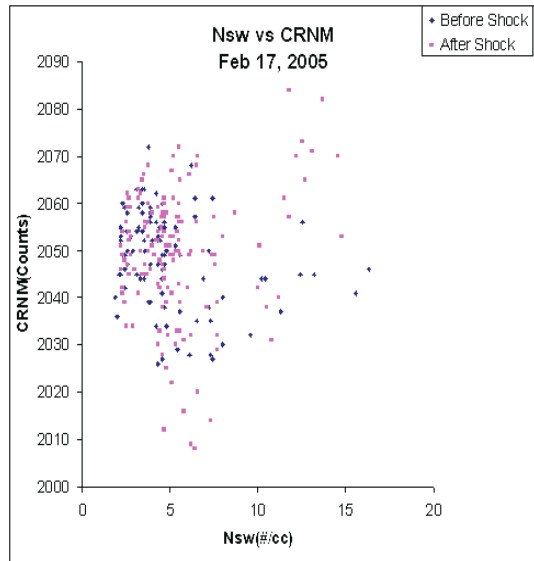


Figure 4(b). Same as Fig. 4(a), but for the diffuse shock event of 17 February 2005. Note that in this case unlike for Fig. 4(a), the normalized MP ratio is not less than 0.8, and the relationship to Nsw is an almost linear one, both before the shock and after the shock. Clearly Nsw is a major parameter in controlling MP. As in the case of Fig. 4(a), the CRNM count shows no relationship with Nsw, thereby ruling it out completely as a controlling parameter.

for ± 5 days from zero time, i.e., the time of IP shock. But due to brevity of space, we are discussing only two cases (one sharp shock type and another diffused shock type event) are discussed in the present paper.

3. Conclusions

- There are definite relationships between the changes in the solar wind/IMF parameters, and the ground-based geomagnetic indices Kp, Dst, CRNM count, and the distance of the dayside magnetopause (MP), notably so after the onset of an IP shock. There is generally clutter of relationship points before the shock.
- These relationships differ for shock events of the sharp type and the diffuse type.
- Solar wind speed Vsw is clearly related to the Kp and Dst indices, and to MP and the CRNM count.
- Solar wind density Nsw clearly controls MP distance, marginally controls Kp and Dst, but does not control CRNM.
- Total B (IMF) is related to changes in Kp and Dst, and clearly controls both MP and CRNM.
- Rapid oscillations in Bz (+ve to -ve and *vice versa*) are well-related to increases in Kp, decreases in Dst, and to decrease in the MP distance and the CRNM count.
- Sunward-directed Bx (+ve) shows association with increased Kp, decreased Dst, and decrease in MP distance and CRNM.
- Downward-directed By (-ve) seems to increase Kp and Dst, but does not show clear relationship with the MP distance and CRNM count.
- Approximately 20 shock events of the year 2005 have been studied, of which the 15 May (sharp shock), and 17 February (diffuse shock) are two representative examples. Through averaging procedures on the 20 shock events, we are establishing quantitative relationships between the changes in the geomagnetic parameters, and the solar wind and IMF parameters. Such quantitative relationships will be invaluable for modelling studies, and for space weather predictions.

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Multipoint Observations of Low Latitude ULF Pc3 Waves in South-East Australia

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Abstract. Geomagnetic pulsations recorded on the ground are the signatures of the integrated signals from the magnetosphere. Pc3 geomagnetic pulsations are quasi-sinusoidal variations in the earth's magnetic field in the period range 10–45 seconds. The magnitude of these pulsations ranges from fraction of a nT (nano Tesla) to several nT. These pulsations can be observed in a number of ways. However, the application of ground-based magnetometer arrays has proven to be one of the most successful methods of studying the spatial structure of hydromagnetic waves in the earth's magnetosphere. The solar wind provides the energy for the earth's magnetospheric processes. Pc3-5 geomagnetic pulsations can be generated either externally or internally with respect to the magnetosphere. The Pc3 studies undertaken in the past have been confined to middle and high latitudes. The spatial and temporal variations observed in Pc3 occurrence are of vital importance because they provide evidence which can be directly related to wave generation mechanisms both inside and external to the magnetosphere. At low latitudes ($L < 3$) wave energy predominates in the Pc3 band and the spatial characteristics of these pulsations have received little attention in the past. An array of four low latitude induction coil magnetometers were established in south-east Australia over a longitudinal range of 17 degrees at $L = 1.8$ to 2.7 for carrying out the study of the effect of the solar wind velocity on these pulsations. Digital dynamic spectra showing Pc3 pulsation activity over a period of about six months have been used to evaluate Pc3 pulsation occurrence. Pc3 occurrence probability at low latitudes has been found to be dominant for the solar wind velocity in the range 400–700 km/s. The results suggest that solar wind controls Pc3 occurrence through a mechanism in which Pc3 wave energy is convected through the magnetosheath and coupled to the standing oscillations of magnetospheric field lines.

Key words. Pc3 magnetic pulsations—Pc3 occurrence—solar wind velocity—magnetospheric physics.

1. Introduction

Observations of geomagnetic pulsations at low latitudes ($L < 3$) indicate that significant hydromagnetic wave energy penetrates deep into the magnetosphere and

the plasmasphere. Statistical studies carried out in the past show that wave energy at low latitudes is primarily in the Pc3 frequency band (Saito 1969; Jacobs 1970; Orr 1973). However, the origin of these waves has not been fully established and it is important to determine whether they are generated within or external to the magnetosphere and to identify their generation mechanism.

It is generally accepted that some of the dayside Pc3 pulsation energy is associated with sources external to the magnetosphere. Statistical studies show that the Pc3 wave period is strongly correlated with the magnitude of the interplanetary magnetic field while the pulsation occurrence rate is dependent on the orientation of the interplanetary magnetic field (Greenstadt *et al.* 1980). The first direct evidence for the propagation of external Pc3-4 wave energy into the magnetosphere had been presented by Greenstadt *et al.* (1983). They had shown that similar wave frequencies were observed simultaneously by ISEE-1 and ISEE-2 spacecraft in the magnetosheath and the outer magnetosphere respectively while lower power was seen within the magnetosphere. In contrast to the external source, waves generated within the magnetosphere must originate from instabilities or free energy sources. Furthermore using the data from GOES-2 satellite, Yumoto *et al.* (1984) have proposed that Pc3-4 wave energy is convected through the magnetosheath to the magnetopause, transmitted deep into the magnetosphere without significant changes in spectra, and then coupled with various hydromagnetic wave modes in the magnetosphere.

There is ample evidence that the solar wind velocity controls some of the properties of Pc3-4 pulsations (Saito 1964; Singer *et al.* 1977). In addition the direction of IMF also plays an important role in controlling these pulsations (Bolshakova and Troistakaya 1968; Takahashi *et al.* 1981). Studies of the joint effect of the solar wind velocity (V_{SW}) and the angle of the interplanetary magnetic field from the Sun–earth line (θ_{XB}) have shown that the amplitude (occurrence) and energy of Pc3-4 pulsations are positively and negatively correlated with V_{SW} and θ_{XB} respectively (Greenstadt *et al.* 1979; Wolfe 1982).

The present study describes the dependence of low latitude Pc3 occurrence on V_{SW} over the period range of March 25 to May 11, 1982. Since IMF data were not available, the Pc3 occurrence dependence on θ_{XB} and the frequency dependence on the IMF magnitude could not be studied.

2. Data and analysis

An array of two-component Pc3-4 induction coil magnetometers were established in south-east Australia to carry out the study (Ansari *et al.* 1985). The array is shown in Fig. 1. The station locations and interstation distances are listed in Tables 1 and 2 respectively.

It can be seen that a geomagnetic longitude range of 17 degrees is covered at $L = 1.8$ and a latitudinal range of 10 degrees over $L = 1.8$ to 2.7. The geomagnetic north–south (X) and east–west (Y) components of Pc3-4 wave signals in the 5–100 mHz were recorded using slow speed analog magnetic tapes employing frequency modulated recording system (Fig. 2). Data from March 25 to September 21, 1982 were digitised in the laboratory with a 5 s sample rate using the recorded time channel pulses and providing a Nyquist frequency of 100 mHz. Digital sonagrams from the X(t) and Y(t) time series of all the data for each recording station over one day (25 h) intervals

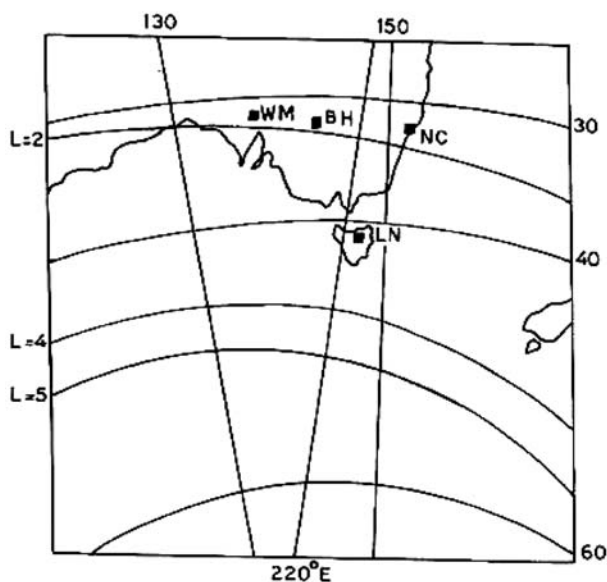


Figure 1. Pc3 recording stations network (WM – Woomera, BH – Broken Hill, NC – Newcastle, LN – Launceston).

Table 1. Co-ordinate details of four Pc3 recording stations.

Station	Geographic		Geomagnetic		L Value
	$\Phi(^{\circ}\text{S})$	$\lambda(^{\circ}\text{E})$	$\Phi(^{\circ}\text{S})$	$\lambda(^{\circ}\text{E})$	
Newcastle (NC)	32.6	151.7	42.0	226.3	1.81
Broken Hill (BH)	32.0	141.5	42.4	214.5	1.81
Woomera (WM)	31.1	136.7	41.7	209.1	1.79
Launceston (LN)	41.7	147.2	52.4	231.1	2.69

Table 2. Interstation distances of the four Pc3 recording stations in Australia.

Station pair	Interstation distance
WM-NC	425 km
BH-NC	1065 km
WM-LN	1485 km
NC-LN	1120 km

were constructed using the Maximum Entropy Method (MEM) on 10-minute subsets overlapping by 5 minutes. A detailed analysis of the digital sonagram characteristics was then carried out and the duration of Pc3 occurrence to the nearest five minutes and the signal frequencies were determined for each station. This information was used to calculate the Pc3 occurrence probability and consequently study the control of V_{SW} on Pc3 pulsations.

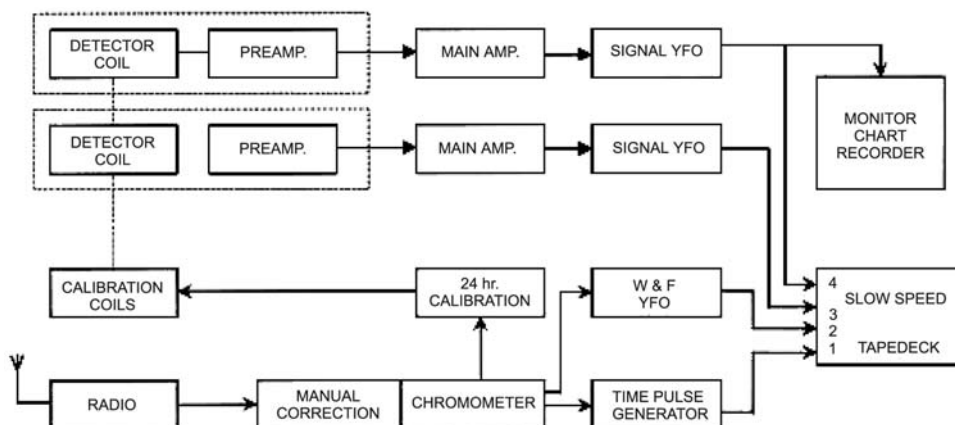


Figure 2. Block diagram of Pc3 recording system.

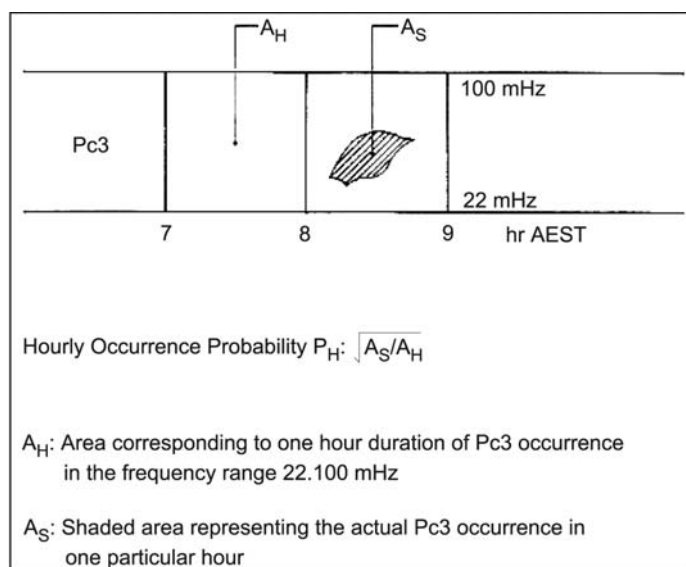


Figure 3. Hourly occurrence probability of Pc3 (after Saito *et al.* 1979).

3. Results

The hourly occurrence probability of Pc3 is defined in Fig. 3 and is based on a method developed by Saito *et al.* (1979). The daytime intervals from 0500 to 1900 h AEST in the digital sonagrams are divided into fourteen hourly domains to eliminate local time effects. The hourly value of Pc3 occurrence probability, P_H , is defined by the formula based on the shaded area A_S and hourly area A_H in the frequency time domain. The solar wind velocities were taken from IMP-8 data (King 1978–1982). Days on which solar wind velocity data were available and days which showed Pc3 activity at one or more stations are shown in Table 3. There were only 16 out of 48 days which provided common solar wind velocity and Pc3 occurrence.

Table 3. Solar wind velocity data and Pc3 occurrence over the time interval 05:00–19:00 h AEST.

Date	V_{SW}	Pc3	Date	V_{SW}	Pc3
25.3.82	–	A	18.4.82	–	A
26.3.82	–	A	19.4.82	–	A
27.3.82	–	A	20.4.82	–	A
28.3.82	–	A	21.4.82	–	A
29.3.82	A	–	22.4.82	A	A
30.3.82	A	A	23.4.82	A	A
31.3.82	A	A	24.4.82	–	A
01.4.82	A	–	25.4.82	A	A
02.4.82	A	–	26.4.82	A	A
03.4.82	A	–	27.4.82	–	–
04.4.82	A	A	28.4.82	A	A
05.4.82	–	A	29.4.82	–	A
06.4.82	–	A	30.4.82	–	A
07.4.82	–	A	01.5.82	–	A
08.4.82	–	–	02.5.82	–	A
09.4.82	–	–	03.5.82	–	A
10.4.82	A	–	04.5.82	A	A
11.4.82	A	–	05.5.82	A	A
12.4.82	A	–	06.5.82	A	–
13.4.82	A	A	07.5.82	A	–
14.4.82	A	A	08.5.82	A	–
15.4.82	A	A	09.5.82	A	–
16.4.82	A	A	10.5.82	A	A
17.4.82	–	A	11.5.82	A	A

Note: A: Solar wind velocity data available or Pc3 occurrence observed.

– : Solar wind velocity data not available or Pc3 occurrence not observed.

The dependence of Pc3 occurrence probability on the solar wind velocity at all the four stations over the 16 days is shown in Fig. 4. The results show a variable pattern with Pc3 occurrence around $V_{SW} = 400$ km/s followed by a gap between 520 and 610 km/s and no occurrence over 700 km/s. Over all velocity ranges, the Pc3 occurrence probability is uniformly high ($\geq 50\%$). The lack of Pc3 activity around $V_{SW} = 600$ km/s may be due to the low occurrence of V_{SW} in this velocity range as illustrated in the V_{SW} distribution shown in Fig. 5.

4. Discussion and conclusion

The dependence of Pc3 occurrence on V_{SW} found in the present study, based on a small data sample is in general agreement with previous studies on the ground (Greenstadt *et al.* 1979, 1980) and in space (Takahashi *et al.* 1981). Greenstadt *et al.* (1979) examined the maximum amplitude of Pc3 for solar wind parameter correlations, while

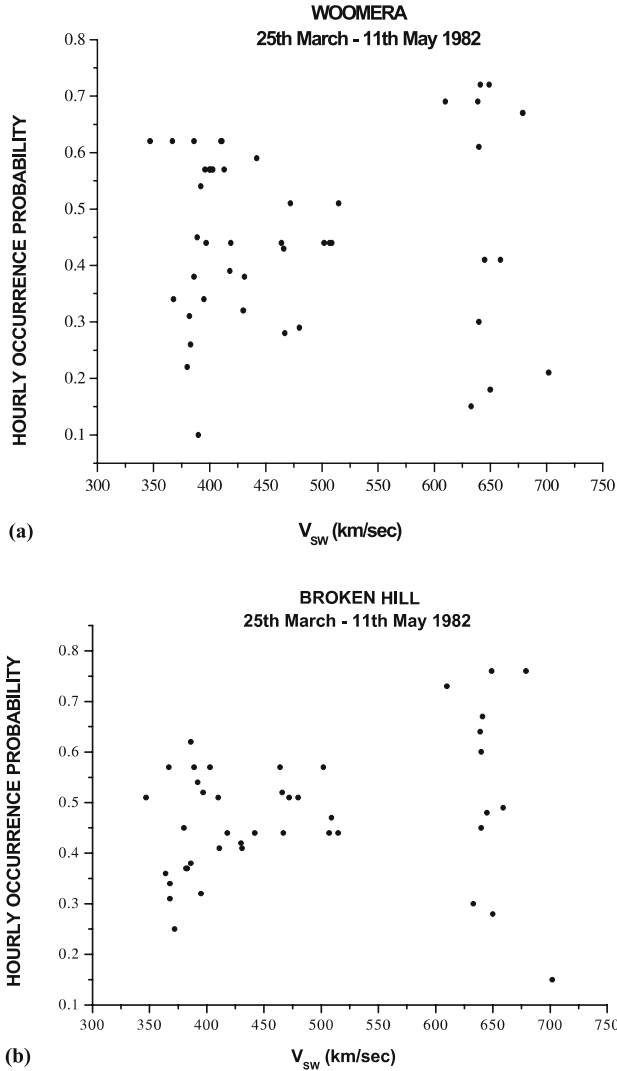


Figure 4. (Continued)

the definition of occurrence of Pc3 pulsations in this study was based on the shape of the spectra. Guglielmi and Potapov (1994) have also reported similar results.

There are two possible locations for the external origin of pulsations, at the magnetopause, and upstream from the magnetopause. Surface waves generated by the K-H instability are important at the magnetopause (Southwood 1968; Boller and Stolov 1973). Upstream from the magnetopause large amplitude waves in the quasi-parallel bow shock are swept back into the magnetosheath and then penetrate the magnetosphere and coupled to the standing oscillations of the magnetospheric field lines (Greenstadt 1972). This coupling may occur at all latitudes (Yumoto and Saito 1983).

The characteristics of the waves excited by the K-H instability are dependent on the length of the field lines and the plasma density at the magnetopause (Chen and

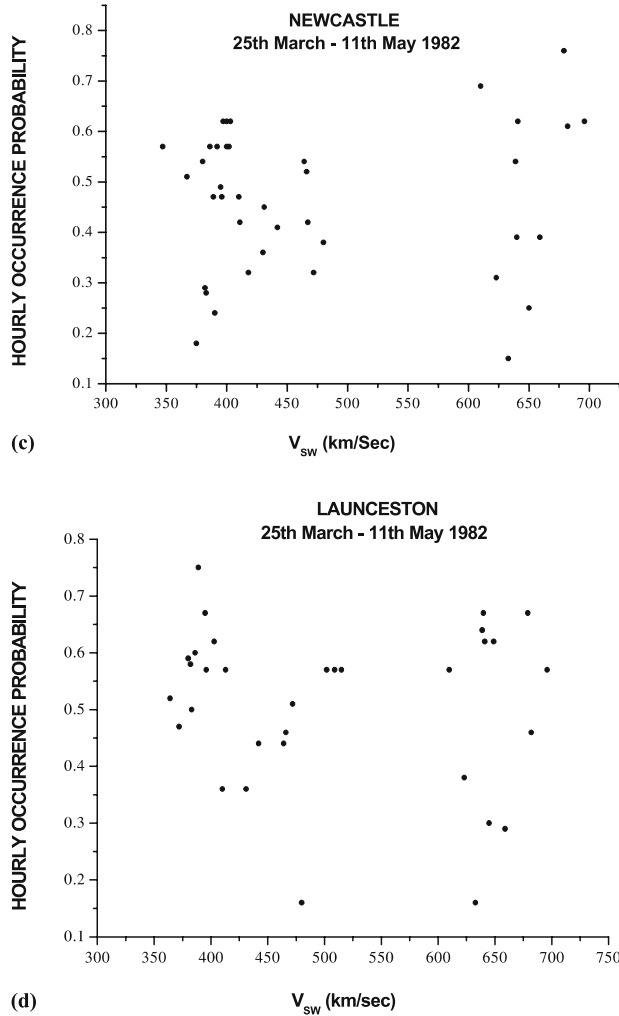


Figure 4. The dependence of Pc3 occurrence probability on the solar wind velocity (V_{SW}) over 16 days in the duration March 25 to May 11, 1982 in the 0500–1900 h AEST interval at (a) Woomera, (b) Broken Hill, (c) Newcastle, and (d) Launceston.

Hasegawa 1974). The association of westward propagation of Pc3 waves with left hand (LH) ellipticity pre-noon and eastward propagation with right hand (RH) ellipticity afternoon for the azimuthal pair of stations (Ansari and Fraser 1985) are consistent with the waves generated at the magnetopause by the K–H instability. The Pc3 noon peaks in occurrence (Ansari and Fraser 1985), however, cannot be explained by this mechanism. It is also difficult to see how the externally excited evanescent surface waves with large damping rates can propagate deep into the magnetosphere and through the plasmopause to couple with the field line resonance at low latitudes. In addition the threshold velocity for the K–H instability involves the angle between the magnetic fields across the magnetopause. Hence on a statistical basis it is likely that the magnetopause is more unstable for higher solar wind velocity, which is consistent with the present results, the contribution of θ_{XB} to Pc3 occurrence cannot be interpreted in the absence of IMF data.

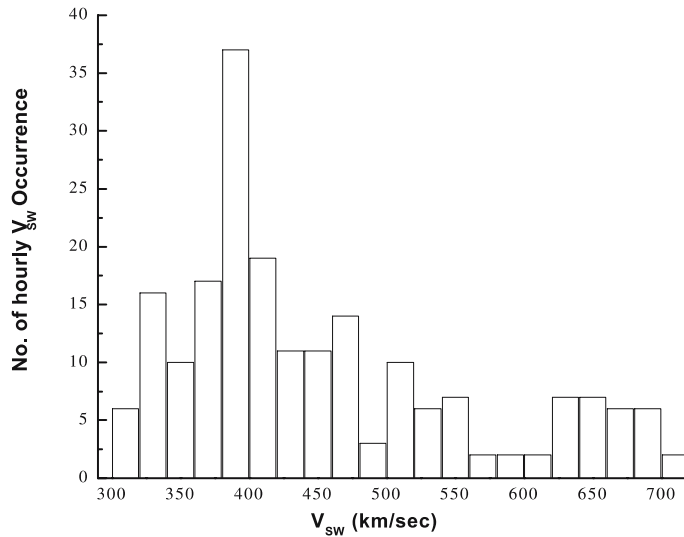


Figure 5. Hourly occurrence of solar wind velocity (V_{SW}) over the period March 25 to May 11, 1982 in the 0500–1900 h AEST interval.

The quasi-parallel shock transition has been found to be highly turbulent (Greenstadt *et al.* 1977). Furthermore, the amplitude of bow-shock-associated waves seems to be dependent on the magnetosonic Mach Number (Formisano *et al.* 1973) and therefore on the solar wind velocity. The association of higher probability of Pc3 occurrence at low latitudes with higher solar wind velocity is therefore more likely to be a consequence of bow-shock associated waves. The results of Yumoto *et al.* (1984) support this mode of Pc3 wave generation and resonance. More recently Chugunova *et al.* (2003), using the data of search coil magnetometers from two Antarctic stations (sub-auroral Pc3), have found that the suggested idea about the possibility of two channels of the penetration of primary upstream turbulence, i.e., via the cusp and via the lobe flanks is statistically feasible. It is however obvious that further studies are needed, which include IMF data, in order to uniquely determine the external and internal contributions to the generation of Pc3 waves observed at low latitudes.

In conclusion it has been demonstrated that the occurrence probability of Pc3 pulsations depends on solar wind velocity with a threshold at about 320 km/s and ranging up to 700 km/s. It is likely that an instability originating from the direct interaction between the solar wind and the magnetosphere is exciting Pc3 pulsations through bow-shock associated waves.

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Energy Dependence of Near-relativistic Electron Spectrum at Geostationary Orbit during the SEP Events of 2005

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Abstract. In view of the renewed interest in the study of energetic particles in the outer radiation belt of the earth, we feel it will be helpful in looking for the energy dependence of the electron energy spectrum at geostationary orbit. This may give us some insight into how we can safeguard geostationary satellites from functional anomalies of the deep dielectric charging type, which are caused by charge accumulation and subsequent discharge of relativistic electrons. In this study we examine whether there is any energy dependence in relativistic electron enhancements at geosynchronous altitudes during solar energetic proton events of 2005.

Key words. SEP events—shocks—geostationary orbit—REDs—REEs.

1. Introduction

The study of charged particle distributions in the earth's radiation belts encompasses as many unknown features as known. One enigmatic area concerns the energetic particles which cause problems in instruments onboard satellites in space; in particular, the Relativistic Electron Dropouts (REDs) and the Relativistic Electron Enhancements (REEs) observed at the geosynchronous altitudes. The sudden drop in the relativistic electron flux often by two orders of magnitude is called RED which is usually preceded by the gradual enhancement in the relativistic electron fluxes, REEs. Certain operational problems on geostationary satellites were attributed to enhanced relativistic electron flux at geostationary altitudes (Baker *et al.* 1987, 1994). The processes which are responsible for electron acceleration in the outer magnetosphere are still not completely known (Rostoker *et al.* 1998; Liu *et al.* 1999). Onsager *et al.* (2002), have discussed the radiation belt electron drop outs with respect to their local time, radial and particle-energy dependence. In this paper we present the energy dependence of REDs and REEs at geostationary orbit for electrons at energies 2 MeV, 0.9 MeV, 0.6 MeV with respect to the solar wind and Interplanetary Magnetic Field (IMF) conditions during the proton events of 2005.

Table 1. List of SEPs observed by both ACE and GEOS-10 during 2005, when the GEOS-10 also observed the relativistic electron drop outs and enhancements at geosynchronous orbit.

Month/year of SEP	Start date	End date
January 2005	15 January	24 January
May 2005	5 May	8 May
May 2005	13 May	16 May
June 2005	16 June	19 June
July 2005	13 July	19 July
July–August 2005	25 July	2 August
August 2005	21 August	26 August
September 2005	7 September	17 September

2. Data presentation

In order to identify the Solar Energetic Particle (SEP) events, we analyzed the fluxes of $E > 10$ MeV and $E > 30$ MeV proton fluxes at both Lagrangian point 1 (L1) and geosynchronous orbit measured by the Advanced Composition Explorer (ACE) and Geosynchronous Earth Observation Satellite-10 (GEOS-10) satellite respectively. These proton fluxes measured by ACE are taken from CDA (Coordinated Data Analysis) web service (<http://cdaweb.gsfc.nasa.gov>) and those measured by GEOS-10 are taken from Space Physics Interactive Data Resource (<http://spidr.ngdc.noaa.gov>). To identify the REE at geosynchronous orbit, we analyzed the data of $E > 0.6$ MeV and $E > 2$ MeV electron fluxes measured by GEOS-10 satellite and differential electron fluxes centered at 0.9 MeV measured by the SOPA instrument on board the LANL 1994, 084 satellite. These electron fluxes are taken from CDA web service (<http://cdaweb.gsfc.nasa.gov>). We are able to identify 8 such events when we see an REE during the SEPs, which are listed in Table 1.

In order to investigate the characteristic solar wind and IMF associated with these events under consideration, we collect hourly solar wind parameters of solar wind number density (N_{sw}), solar wind speed (V_{sw}), solar wind pressure (P_{sw}), average component of IMF (B_{av}) and z-component of IMF (B_z) measured by ACE Satellite. These data are taken from CDA web service (<http://cdaweb.gsfc.nasa.gov>).

In this paper we present a representative event of January 2005, for all the events under consideration. We discuss the common features observed in all the events. Figure 1 shows the proton fluxes observed by both ACE and GEOS-10 satellites, confirming the SEP during this period. In Fig. 2, we present the characteristic solar wind and IMF behaviour and the geosynchronous electron response. In Fig. 2 (from top to bottom) we present V_{sw} , N_{sw} , P_{sw} , B_{av} , B_z , and electron fluxes of > 2 MeV, 0.9 MeV and > 0.6 MeV respectively.

3. Summary of work

In this section we discuss the observations and our calculations. For the representative event from Fig. 1, we can say that there is SEP observed by both ACE at L1 point and GEOS-10 at geosynchronous orbit. From the bottom 3 panels of Fig. 2, it is evident that the REEs are observed by the GEOS-10 and LANL satellites. The scale is in terms of

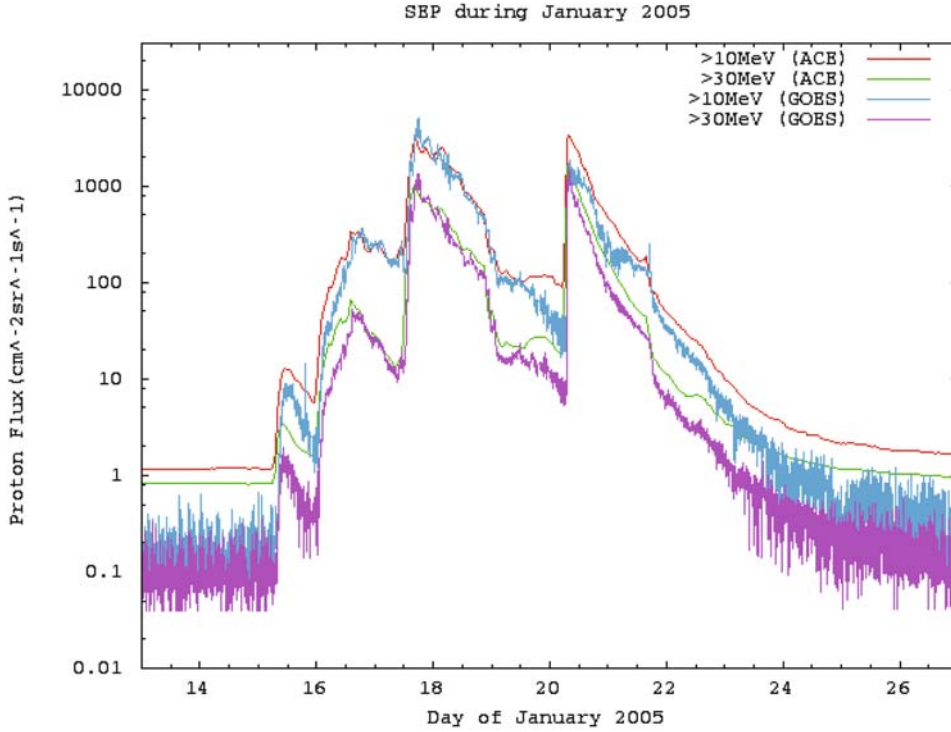


Figure 1. The SEP during January 2005 observed by both ACE at Lagrangian point 1 and GEOS-10 at geosynchronous orbit.

particle flux units. From both Figs. 1 and 2, it is clear that there are 2 REEs during this event. In both cases the amount of enhancements and drop outs in the electron fluxes at geosynchronous orbit is different. The drop out and enhancement during the late hours of 16 January is less intense than 21 January. This is because of the difference in the solar wind and IMF behaviour in both the events. Many of the previous workers (Williams 1966; Paulikas and Blake 1976; Li *et al.* 2001a, 2001b; Kim *et al.* 2006) who concentrated in the acceleration and loss mechanisms of relativistic electrons in the magnetosphere, showed how the solar wind and IMF controls the relativistic electron changes in the magnetosphere of earth. The solar wind and interplanetary drivers during the 17 January case are less pronounced than in the 21 January case. In particular Psw during 17 January is very less, which we believe is the main cause for the intense drop out and enhancement seen in the 21 January case. In this study we tried to look at the enhancements and drops of the electron fluxes during the SEPs.

We have calculated the amount of drop and amount of enhancements during the REDs and the REEs respectively. To calculate the amount of drop during the REDs we took the ratio of the maximum and minimum flux values before and after the occurrence of REDs. This represents the amount of drop in the electroflux during RED. Similarly, we have taken the ratio of maximum to minimum values of electron flux after and before the REEs respectively. This represents the amount of enhancement in the electron flux during REE. In the representative case we found that the amount of drop and enhancements of > 2 MeV electrons is $4.2e + 5$ and $4.6e + 4$ respectively.

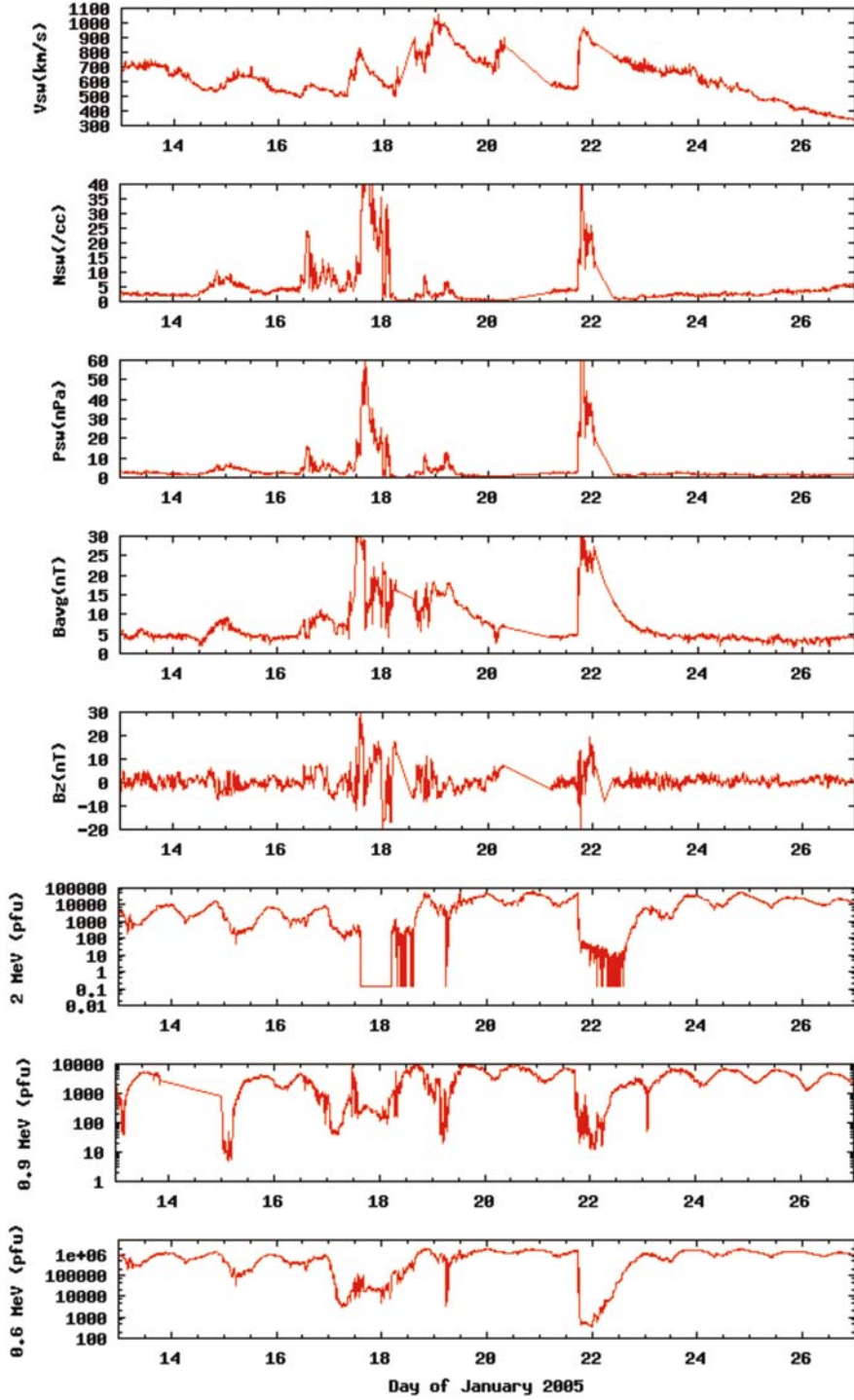


Figure 2. The solar wind and IMF behaviour is shown, along with the geosynchronous electron flux response. From top to bottom, Vsw, Nsw, Psw, Bav of IMF, Bz of IMF, 2 MeV electron flux, 0.9 MeV electron flux and 0.6 MeV electron flux respectively.

Similarly, the drop and enhancements observed by 0.9 MeV are $1e + 3$ and $4.8e + 2$ respectively and for > 0.6 MeV, it is $4e + 3$ and $3e + 2$ respectively. The amount of drop and enhancements are more for > 2 MeV electron than for > 0.6 MeV and 0.9 MeV electron fluxes. This kind of behaviour is observed in all the events under consideration.

4. Conclusions

The above-mentioned observations made us think that the solar and interplanetary disturbances which are responsible for the REEs at geosynchronous altitudes, may be driving different kinds of acceleration and loss mechanisms that are acting differently on electrons with different energies or these acceleration and loss mechanisms are acting predominantly on some energies than on the others. But from our observations we are convinced that there is some kind of energy dependence in the way the acceleration and loss processes at geosynchronous altitudes are working.

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Next Generation UV Coronagraph Instrumentation for Solar Cycle-24

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Abstract. Ultraviolet coronagraph observations of the extended solar corona (defined here as 1.5 to 10 solar radii from Sun-center) have become a powerful tool for obtaining detailed empirical descriptions of coronal holes, streamers, and coronal mass ejections. The empirical models resulting from ultraviolet coronagraph observations provide the constraints needed to test and guide theoretical models aimed at determining the physical processes that control solar wind acceleration, CME heating and acceleration, and solar energetic particle (SEP) acceleration. Measurements to date from sounding rockets, the shuttle deployed Spartan 201 satellite and the Solar and Heliospheric Observatory (SOHO) have utilized high resolution spectroscopy over a very limited instantaneous field of view. New concepts for next generation instrumentation include imaging ultraviolet spectro-coronagraphs and large aperture ultraviolet coronagraph spectrometers. An imaging instrument would be the first to obtain absolute spectral line intensities of the extended corona over a wide field of view. Such images would provide the absolute intensities of spectral lines that can be used to determine densities and outflow velocities of specific coronal ions. Measurements from several charge states of a given element will allow electron temperatures to be determined. These measurements combined with observations of H I Ly α provide absolute chemical abundances (relative to hydrogen) for observed elements. Ultraviolet imaging would be highly complementary to a large-aperture ultraviolet coronagraph spectrometer designed for high spectral resolution observations over a small instantaneous field of view. The images would be used to select targets for more detailed spectroscopic studies with the large aperture UV coronagraph spectrometer and to provide time dependent empirical descriptions of the regions surrounding the narrow instantaneous field of view of the large aperture instrument. Descriptions of both the imaging ultraviolet spectro-coronagraph and the large aperture ultraviolet coronagraph spectrometer are provided. Recommended co-observing instruments are described.

Key words. Solar instrumentation—solar spectroscopy—solar corona—solar wind—coronal mass ejections.

1. Introduction

The first ultraviolet spectroscopy of the extended solar corona in the absence of a natural solar eclipse was carried out in April 1979 when a rocket-borne ultraviolet coronagraph spectrometer measured absolute intensities and spectral line profiles of H I Ly α in a streamer and a polar region at heliographic heights from 1.5 to 3.0 solar radii ($R_{\odot}$) from Sun-center (Kohl *et al.* 1980). Since that time, ultraviolet spectroscopic observations of the extended corona have been carried out on two sounding rocket flights, four shuttle deployed Spartan 201 flights and during the 12-year history of the Ultraviolet Coronagraph Spectrometer (UVCS) instrument on the Solar and Heliospheric Observatory (SOHO) satellite (Kohl *et al.* 1995). The overall scientific goal of all of these observations has been to obtain detailed empirical descriptions of coronal structures and to use these descriptions together with theoretical models to identify and understand the physical processes that release, transport, and deposit energy in the corona and produce the solar wind, coronal mass ejections (CMEs), and solar energetic particles. For a review of ultraviolet spectroscopy of the extended solar corona, see Kohl *et al.* (2006).

Although considerable progress has been made, new instrumentation will be needed to build on the discoveries of the earlier observations. A combination of imaging and spectroscopic diagnostic techniques are needed to provide more complete empirical descriptions capable of testing and guiding the development of next generation theoretical models. Imaging uniquely provides the global morphology and dynamics of the densities, flow speeds and chemical abundances of the often complex and filamentary structures of coronal plasmas. Spectroscopy uniquely provides ion temperatures, anisotropic velocity distributions, electron non-Maxwellian velocity distributions and suprathermal seed particle populations needed for CME shock acceleration of solar energetic particles. By comparing the empirical descriptions to the predictions of theoretical models, we can discard models that do not agree with the measurements and build the case for the model and associated process(es) that agree with the measurements.

2. Primary results from Ultraviolet Coronagraph Spectrometers

In the case of the solar wind, it was discovered from UVCS/SOHO and earlier instruments that the fast solar wind becomes supersonic much closer to the Sun than previously believed. In coronal holes, heavy ions (e.g., O⁵⁺) both flow faster and are heated hundreds of times more strongly than protons and electrons, and they have anisotropic temperatures. Extended heating and acceleration remain strong upto and beyond the largest heights observable with UVCS. These observations have rekindled theoretical efforts to understand coronal heating, concentrating on cyclotron resonance of kHz-frequency waves in the collisionless extended corona where the primary solar wind acceleration occurs. Slow wind from bright quiescent equatorial streamers at solar minimum was observed to flow along open-field “edges” and exhibits similar high ion temperatures and anisotropies as coronal holes (but at larger heights). It remains to determine what fraction of the slow wind comes from these boundary regions and what is produced by transient eruptions.

Even though UVCS/SOHO has made significant advances, we still do not understand the physical processes that heat and accelerate the entire plasma (protons, electrons, helium, and minor ions). Our understanding of ion cyclotron resonance is based

essentially on just one ion, O^{5+} . There is still controversy about whether the fast solar wind occurs primarily in dense polar plumes or in low-density inter-plume plasma. We also do not know how and where the various components of the variable slow solar wind are produced. UVCS/SOHO has shown that answering these questions is possible, but it cannot make the observations needed to fully characterize the plasma.

In the case of CMEs, UVCS observations together with observations by other SOHO instruments have demonstrated the power of combining images and spectroscopy to reveal physics that no one instrument could discern alone. There have been several key results. The first spectroscopic determination of CME thermal energy was surprisingly found to be as large as the outflow kinetic energy. UVCS observations were used to make the first determinations of CME helical rotation rates and handedness in the extended corona, which are key features of theoretical CME models. Measurements were made of the properties of shock fronts and reconnection current sheets. UVCS data have been used to characterize sites of solar energetic particle (SEP) production (e.g., CME shocks) for specific events, needed to test and guide SEP acceleration/transport models.

Even though UVCS/SOHO has made significant advances, we still do not have a comprehensive empirical model of a CME that would provide the key plasma parameters as a function of time, and the information needed to identify the physical processes that heat and accelerate CMEs. Many questions about CME evolution remain unanswered. For example: is the bright “leading edge” of the CME an arcade of active region magnetic loops, is it ambient coronal gas compressed by a shock, or is it a “bubble” of coronal plasma created by reconnection as the CME expands? What is the role of magnetic reconnection in both the eruption and relaxation of the coronal magnetic field in CMEs? UVCS has shown that answering these questions is possible, but it cannot make the observations needed to fully characterize the plasma.

3. An Ultraviolet Imaging Spectro-Coronagraph

An Ultraviolet Imaging Spectro-Coronagraph payload can consist of one or more telescopes such as that depicted in Fig. 1. Light from the solar disk enters the instrument through an entrance aperture formed by a circular external occulter and an outer aperture stop. This light passes by a series of baffles and the primary telescope mirror, and into a light trap where it is attenuated and discarded. The light from the extended corona passes through the same aperture, past the baffles and onto the telescope mirror. The external occulter blocks the light from the solar disk that would otherwise impinge on the telescope. The telescope mirror forms an image of the extended corona at the prime focus. Light forming the image passes through a field stop at the prime focus, past the internal occulter and onto a concave diffraction grating located near the entrance aperture. The mirror forms an image of the external occulter on the internal occulter, which blocks diffracted disk light from the external occulter. The grating disperses the coronal light and focuses it through a hole in the telescope mirror and onto the face of an array detector. In this way, overlapped images of the extended corona in selected spectral lines are detected. Several techniques can be used to reconstruct images of the extended corona in individual spectral lines from the detected overlapped images. Shadow sensors located in the Sunlight trap are used to determine the alignment of the occulting system with the Sun.

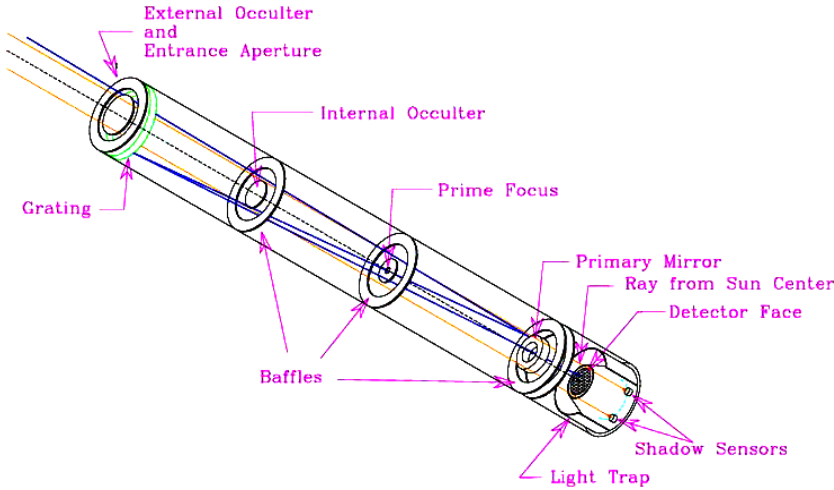


Figure 1. Isometric optical diagram of a single ultraviolet spectro-coronagraph telescope.

An imaging spectro-coronagraph payload in geostationary orbit would be expected to observe approximately 1000 CMEs per year. A relatively small instrument could provide the needed spatial and spectral resolution in a set of spectral lines covering a broad temperature range. Spectral lines observable with such an instrument over a range of heliographic heights from 1.5 to 3.5 $R_{\odot}$ include H I 121.6 nm, OVI 103.2 and 103.7 nm, C III 97.7, Fe XVIII 97.5 nm, He II 30.4 nm, Si IX 29.6 nm, Si X 27.7 nm, Si XII 49.9 and 521 nm and Mg X 61.0 and 62.5 nm.

Ultraviolet spectroscopy of the extended corona has been limited to observations of narrow strips of the corona. An imaging spectro-coronagraph could measure, simultaneously, the absolute intensities of several spectral lines over a large portion of the corona. An imaging spectro-coronagraph could measure the helium abundance and He^+ flow speed in CMEs, coronal holes and coronal streamers for the first time. Helium is the second greatest contributor to the mass and momentum of CMEs and the solar wind, and yet we know little about its abundance and flow speeds in the corona. Measured densities, abundances and outflow velocities of a broad range of ions are expected to provide breakthroughs in identifying how CMEs are formed, and how they evolve.

4. An advanced Ultraviolet Coronagraph Spectrometer

The basic optical design concepts for next generation ultraviolet coronagraph spectrometers are similar to that of UVCS/SOHO except that the external occulter is placed much further from the telescope mirror in order to achieve a much larger effective area and higher diffraction limited resolution in the radial direction. The basic optical designs of all ultraviolet coronagraph spectrometers flown to date are similar to the UVCS/SOHO design shown in Fig. 2. It consists of one or more articulated telescope mirrors that are placed in the shadow created by a linear external occulter, which forms one side of a rectangular entrance aperture. The mirror is used to image a portion of the extended solar corona onto the entrance slit of an ultraviolet spectrometer.

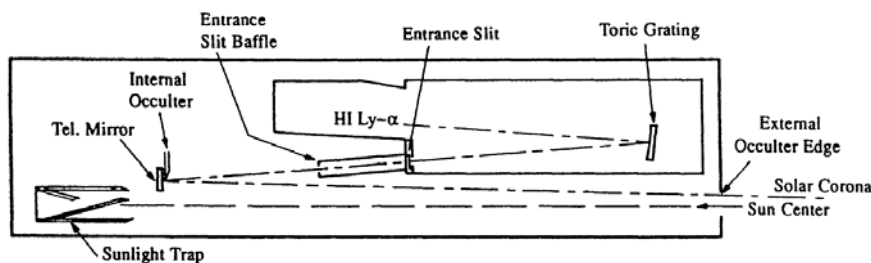


Figure 2. Optical layout of the UVCS/SOHO H I Ly α channel (from Kohl *et al.* 1995).

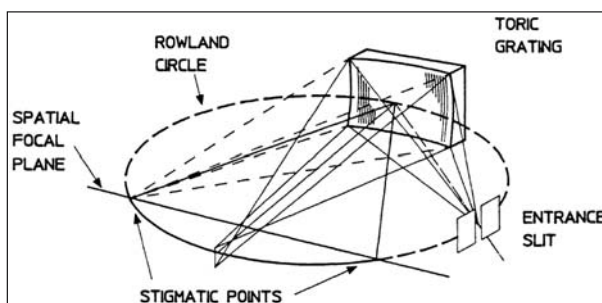


Figure 3. Isometric diagram of a toric grating spectrograph similar to that of UVCS/SOHO (from Kohl *et al.* 1995).

A linear internal occulter placed near the telescope mirror, intercepts and removes solar disk light that is diffracted towards the mirror and would otherwise be specularly reflected by the mirror onto the spectrometer entrance slit. This externally and internally occulted telescope design provides the stray light suppression at the wavelengths of interest that is needed to observe the relatively faint ultraviolet extended corona. Stray solar disk light at wavelengths outside the wavelengths of interest is suppressed by a combination of the occulting system, the optical coatings, the dispersive action of the diffraction grating and the wavelength dependent sensitivity of the detector. A Sunlight trap intercepts and absorbs solar disk light that passes through the entrance aperture, past a series of baffles, and past the telescope mirror. The spectrometer entrance slit is shielded from the illuminated surfaces of the trap by an entrance slit baffle and the telescope mirror. Each channel of the UVCS/SOHO spectrometer has an entrance slit, a toric grating and an array detector mounted in a Johnson–Onaka (Johnson 1957) configuration (see Fig. 3). The two detectors for the two ultraviolet channels are two-dimensional photon counting, centroiding, microchannel plate sensors with electronic readout (Siegmund *et al.* 1994).

The spatial resolution in the radial direction and the effective area are extremely limited in ultraviolet coronagraph spectrometers such as UVCS/SOHO. For example, the 1.8 m separation between its external occulter and its telescope mirror, together with the need to further occult with the internal occulter, results in an unvignetted telescope width of 0.8 mm when observing at $1.5 R_{\odot}$. The limitation can be greatly improved with a remote external occulter supported by an extendable boom. Space qualified booms of 13 m and longer are available that can provide over 14 m separation

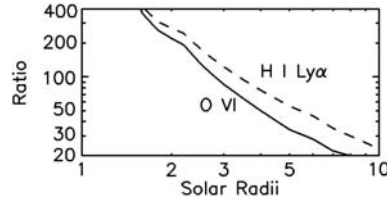


Figure 4. Ratio of the efficiencies of an advanced large-aperture ultraviolet coronagraph spectrometer to those of UVCS/SOHO for observations with the same spatial and spectral resolutions at H I Ly α and O VI 103.2 nm (from Kohl *et al.* 2006).

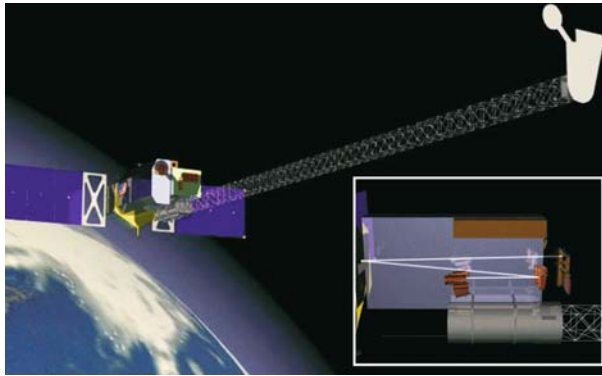


Figure 5. Spacecraft concept for advanced large-aperture coronagraphs. The inset is a diagram of the telescope/spectrometer unit of a next generation ultraviolet coronagraph spectrometer (from Kohl *et al.* 2006).

between the external occulter and the telescope mirror. This arrangement provides an unvignetted mirror width of 27 mm for observations at $1.5 R_{\odot}$. Much larger unvignetted widths are provided for observations at larger heights. If telescope mirrors that are about the same size as those of UVCS/SOHO are used (width = 75 mm), then they are filled when observing at $2.1 R_{\odot}$ or higher. This advantage together with improvements in mirror reflection coatings and detector sensitivity provide the gain in effective area shown in Fig. 4. An illustration of a spacecraft concept containing a large aperture ultraviolet coronagraph spectrometer and a co-observing visible light coronagraph is shown in Fig. 5. The inset shows a diagram of the telescope/spectrometer unit of the ultraviolet coronagraph spectrometer whose entrance aperture lies in the shadow of the external occulter.

Each channel of the spectrometer is optimized for spectroscopic measurements in a specific wavelength band. The ideal spectrometer for each channel would have a large radiometric throughput for the wavelengths of interest, stigmatic imagery over the entrance slit length for the primary spectral lines of that channel, good image quality for other spectral lines of interest, and a dispersion and detector pixel size that is compatible with the required spectral and spatial resolution. To accomplish the scientific objectives, it is desirable to have three channels. One channel would cover the 26–37 nm wavelength range and be capable of observing He II 30.4 nm. Another would cover 75–150 nm in first order and 49–75 nm in second order. This channel would be

capable of observing a large number of spectral lines including H I 121.6 nm, OVI 103.2 and 103.7 nm, Fe XVIII 97.5 nm and many more. A separate optical path in that channel could provide polarimetry measurements that could be used to explore the possibility of determining coronal magnetic fields with the Hanle effect. A third channel could have a crossed dispersion spectrometer for observing the electron scattered wing of the H I Ly α spectral line in order to determine non-Maxwellian velocity distributions and temperatures of coronal electrons. The crossed dispersion spectrograph is needed to reduce the scattered light from the diffraction grating by about five orders of magnitude with respect to the resonantly scattered core of the line.

5. Desirable co-observing instruments

It is highly desirable to have a co-observing visible light coronagraph–polarimeter in any science payload that includes an ultraviolet coronagraph spectrometer. The visible light instrument should cover the field of view of the ultraviolet instrument, and ideally would provide two dimensional imaging of the corona upto at least six solar radii. The visible light instrument provides the electron density over its field of view and it provides the time dependent distribution of the electron density. This information is needed to complement the many plasma parameters of coronal structures that can be determined from the ultraviolet instrument observations. The combination of the two datasets can be used to build a time dependent empirical model of the observed coronal structures. It is highly desirable for the visible light coronagraph to also have a spectroscopic capability, which could obtain spectral line profiles near the limb with high spatial and temporal resolution.

Observations of dynamic phenomena such as flare/CME events would also benefit from X-ray observations of the flaring region. Ideally, the X-ray instrument could provide both spectra and imaging.

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High Resolution Observations using Adaptive Optics: Achievements and Future Needs

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Abstract. Over the last few years, several interesting observations were obtained with the help of solar Adaptive Optics (AO). In this paper, few observations made using the solar AO are enlightened and briefly discussed. A list of disadvantages with the current AO system are presented. With telescopes larger than 1.5 m expected during the next decade, there is a need to develop the existing AO technologies for large aperture telescopes. Some aspects of this development are highlighted. Finally, the recent AO developments in India are also presented.

Key words. Small-scale structures—Adaptive Optics—Multi-conjugate Adaptive Optics—high spatial resolution.

1. Introduction

The Sun provides a unique opportunity for a greater understanding of the physical processes happening in stellar atmospheres. The solar atmosphere is structured to very small scales which are dynamic in nature. Understanding the nature of these small-scale structures and dynamics may provide important clues to some puzzling and unanswered questions. It is now very well realised that small-scale dynamics can influence large-scale structures (DeRosa 2005). In hydrodynamic condition, two important scales determine the structuring of the solar atmosphere: (i) pressure scale height, and (ii) photon mean free path. Both these scales correspond to a value of 70 km (i.e., about 0.1") at the photosphere. However, smaller magnetic structures are seen in realistic numerical magneto-hydrodynamic simulations (Stein & Nordlund 2006; Schussler & Vogler 2006). Grid sizes used in these simulations are smaller than 0.1" and future simulations are aimed to achieve a grid size of 0.02" or better.

A few years earlier, resolving structures on the Sun to better than an arcsecond for longer durations were not possible even at good sites due to the earth's atmosphere. The scenario changed with the advancement of the Adaptive Optics (AO) technology. Hence, modern telescopes thrive on improving the spatial, spectral, and temporal resolutions using an AO system. It may be stated that most of the recent large solar

[†]NSO is operated by the Association of Universities for Research in Astronomy (AURA) under co-operative agreement with National Science Foundation (NSF).

Table 1. A list of successful AO systems.

AO	Actuator	Sub-aperture	Camera	Frame rate	Hardware	DM	Ref.	First light
76 cm DST Lockheed	57	19	—	Analog	—	—	Acton 1992	1986
76 cm DST LOAO	97	24	—	< 1.6 kHz	24 DSPs	Xenitics Inc.	Rimmele 2000	1998
48 cm SVST	19	19	Dalsa CA-D6	955 Hz	566 MHz Alpha	AOPTIX Tech. Inc.	Scharmer 2000	1999
76 cm DST HOAO	97	76	Custom built CMOS	2.5 kHz	40 DSPs	Xenitics Inc.	Rimmele 2003	2002
70 cm VTT KAOS	35	36	Dalsa CA-D6	955 Hz	8X900 MHz Sun	Laplacian Optics	von der Luhe 2003	2002
1.5 m McMath Low cost	37	120–200	Dalsa CA-D6	955 Hz	1 GHz Pentium	Okotech	Keller 2003	2002
97 cm SST	37	37	Dalsa CA-D6	955 Hz	1.45 GHz Athlon	AOPTIX Tech. Inc.	Scharmer 2003	2003
65 cm BBSO HOAO	97	76	Custom built CMOS	2.5 kHz	40 DSPs	Xenitics Inc.	Denker 2007	2004

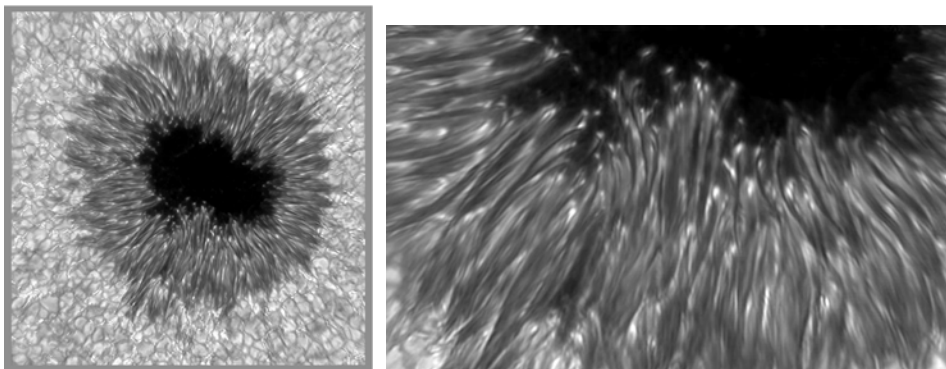


Figure 1. Speckle reconstructed sunspot observed at the DST using the high-order AO (HOO) system. Courtesy: F. Woeger, T. Rimmele, M. Komsa, C. Berst, S. Fletcher & S. Hegwer: NSO/AURA/NSF.

ground-based telescope (aperture ≥ 50 cm) designs have an integrated AO system. The realisation of the day-time AO technology lagged behind the night-time due to the difficulties caused by the extended nature of the source as well as the poor seeing conditions prevailing during the day-time (Rimmele 2004; Keller 2005).

The first solar AO system (Acton 1992, 1993) was developed at Lockheed and tested at the Dunn Solar Telescope (DST). This system worked with high contrast features like solar pores and hence severely limited the scientific use. The development of the correlating Shack–Hartmann wavefront sensor, for the low-order AO system at DST (Rimmele 2000), has changed the perspective of the solar AO. All solar AO systems currently in operation are based on the correlating Shack–Hartmann sensor. Following Keller (2005), Table 1 lists the successful AO systems at different solar observatories around the globe. Figure 1 shows an example sunspot image observed under good seeing conditions using a high-order AO (HOO) system operated at DST. A small portion of the bottom penumbral region is blown up to show the high spatial resolution achieved with this observation.

2. Science achievements

There are innumerable science observations that made use of the AO. Only few observations are briefly discussed in this paper due to page restrictions. The readers are requested to go through the cited references for details.

2.1 High resolution imaging

Narrow band G-band images are the first images from most of the AO systems. This is due to the presence of high contrast small-scale features at this wavelength band. However, watching the AO corrected images in the broad-band video camera attached to the AO system, is always a pleasure for observers. Contrast of these images are usually enhanced using post-processing techniques: speckle reconstruction, phase-diversity, or long-exposure point spread deconvolution with or without frame selection. Each of these techniques has its own advantages and disadvantages.

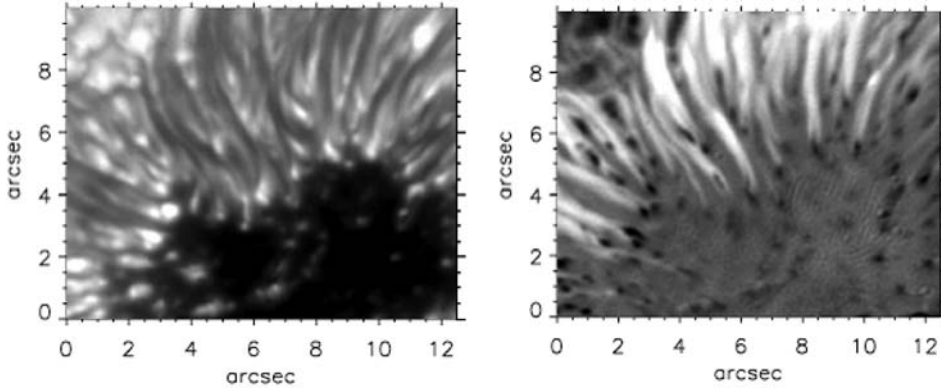


Figure 2. Intensity (right) and velocity (left) image of an observed sunspot penumbral filaments. Dark (bright) in the velocity image represents an upflow (downflow).

One of the best high spatial resolution observation achieved during the initial phase of the AO development is the detection of dark cores in the penumbral filaments (Scharmer *et al.* 2002) using the Swedish Solar Telescope (SST). The observations are obtained using a low order AO system along with real-time frame selection and subsequent image restoration technique. Observations done with HOAO at DST confirmed the existence of dark cored penumbral filaments (see Fig. 1). These observations show that the dark cores are unresolved with a cross-section of about 90 km or smaller and their length even as long as 1000 km. They can live for at least an hour. Apart from the dark cores in the penumbral filaments, other thin dark structures, named as ‘canals’ and ‘hairs’ are also observed. There are other interesting observations from the SST over the last few years. Few examples are: the 3D structures of small-scale fields (Lites *et al.* 2004), leakage of photospheric oscillations and flows to form spicules (de Pontieu *et al.* 2004), multi-line spectroscopic study of the dark cored filaments (Bellot Rubio *et al.* 2005), and the detailed study of the inclination of magnetic fields and flows in the sunspot penumbrae (Langhans *et al.* 2005).

High resolution time sequence observation of velocity fields in the penumbral regions (Rimmele & Marino 2006a) has allowed one to quantitatively compare different penumbral models. The time sequence clearly shows that the upflows (in the footpoints of penumbral dark filaments) and horizontal flow (the evershed flow) move around and evolve as a unit, indicating that they are part of the same feature. These observations also produce strong evidence that the penumbral grains are the inner footpoints of the evershed flows where the hot upflow occurs. Full vector magnetic field observations along with such velocity observations may well be able to quantitatively differentiate and refine the few existing models of the penumbra. Such kind of observations can be expected in the near future. Penumbral asymmetries are also studied in detail with observations obtained using AO (Tritschler *et al.* 2004; Soltau *et al.* 2005). These recent high spatial resolution penumbral observations have already started the debate on the suitability of the existing penumbral models (Spruit & Scharmer 2006; Thomas *et al.* 2006).

Success of the AO has initiated the deployment of several new backend instruments over the past few years. The Interferometric Bidimensional Spectrometer (IBIS) is one such instrument deployed at the DST in 2003. This instrument produces high

spatial, spectral, and temporal resolution observations both in the photosphere and in chromosphere (Cavallini & Reardon 2006). Analysing the time sequence observations obtained with this instrument, Vecchio *et al.* (2007) have concluded that waves with frequencies above the acoustic cut-off propagate to upper layers only in restricted areas of the quiet Sun. Their results support that the network magnetic elements can channel low-frequency photospheric oscillations into chromosphere, thus providing a way to input mechanical energy in the upper layers.

2.2 High resolution spectroscopy

The real test for any solar AO comes from the spectroscopic and spectro-polarimetric observations due to the long exposure times and the time required for scanning the region of interest. Moreover, the raster scan images obtained with a spectrograph does not undergo any special processing (like the speckle reconstruction or phase-diversity). In most cases, the solar AO helps in keeping consistent image quality during the scanning time of the spectrograph. The scanning time can be as large as one hour depending on the region of interest. The initial observations using the low-order AO corrected images along with the Advanced Stokes Polarimeter (ASP) has revealed small-scale needle-like convective structures around a pore (Sankarasubramanian & Rimmele 2003). The radius of the ring-like upflows seen around the small-pore matched quite well with the pattern observed in the magneto-hydrodynamic numerical simulations (Steiner 1998). Figure 3 shows a raster image obtained by scanning the Diffraction Limited Spectro-Polarimeter (DLSP; Sankarasubramanian *et al.* 2004) operating at the DST along with the HOAO. This raster image is an example of the consistent performance of the HOAO for long (45-minute for this case) duration observations.

Regular observations of diffraction-limited vector magnetic field are feasible using the HOAO and DLSP. These high spatial resolution vector magnetic field measurements are started to produce quantitative information about small-scale magnetic elements, like umbral dots (Socas-Navarro *et al.* 2004a; Sankarasubramanian *et al.* 2004b), small-scale magnetic elements in and around active regions (Sankarasubramanian & Hagenaar 2007) and quiet Sun magnetic fields (Lites & Socas-Navarro 2004). There are other similar spectroscopic instruments developed at the AO corrected image planes, examples are the POLIS and TIP (Schmidt *et al.* 2003; Martinez Pillet *et al.* 1999). The development of the AO has also helped in obtaining very high resolution spectroscopic observations in the near and far infra-red regions (Keller *et al.* 2003; Lin, private communication).

3. AO developments in India

The development of AO technology was started very recently in India. A low-cost AO system along with a Shack–Hartmann wavefront analyser was developed and tested in the laboratory (Ganesan *et al.* 2005). A 10×10 sub-aperture for sampling the wavefront and a 37-channel MMDM from OKO Tech was used in this system. This system cannot be used for solar observations due to the slow read-out speed of the camera (25 fps).

Udaipur Solar Observatory (USO) is involved in developing an Adaptive Optics system for their 50 cm aperture Multi-Application Solar Telescope (MAST; Venkatakrishnan 2007). A laboratory set-up has been developed for testing the whole

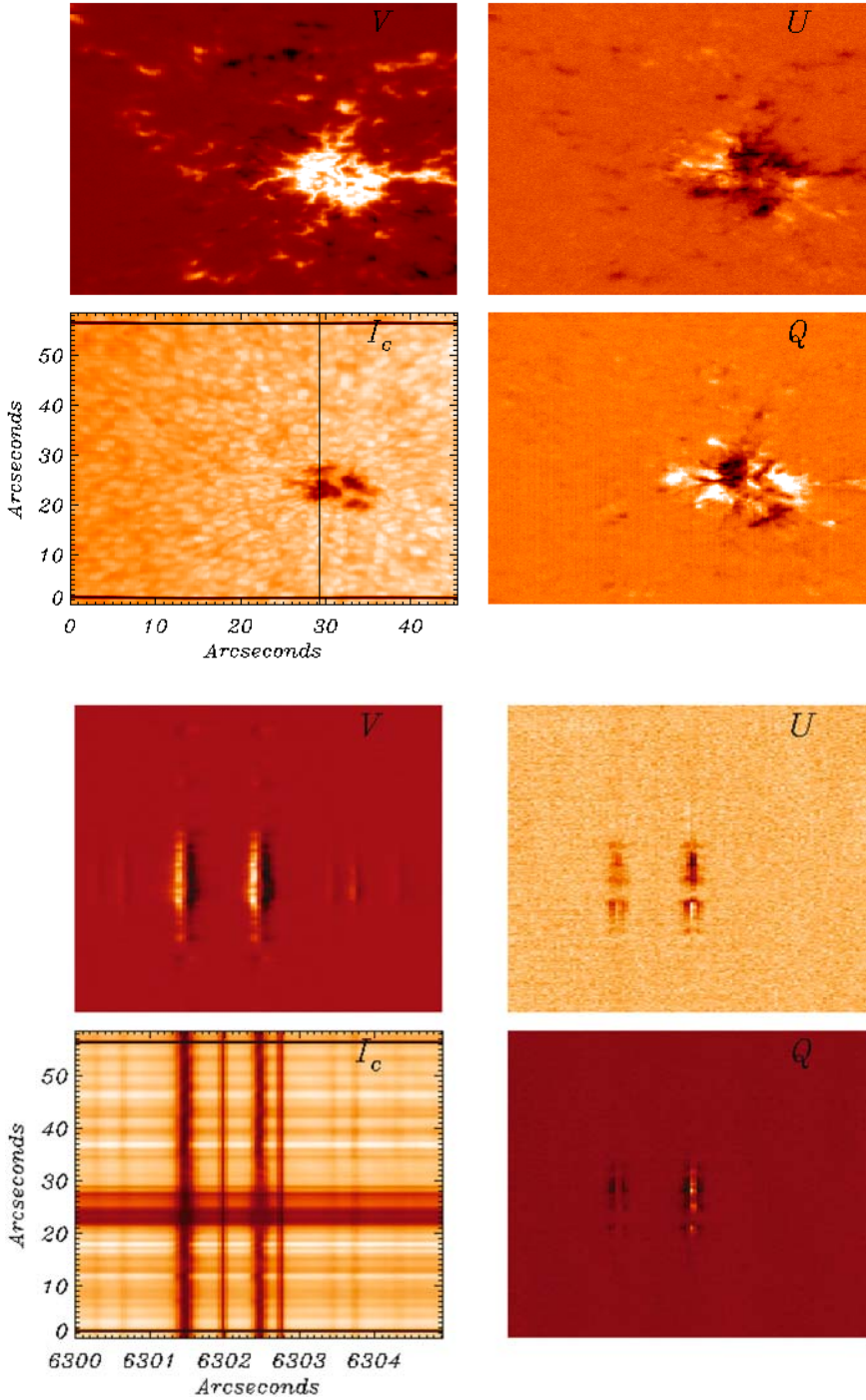


Figure 3. Raster image of the continuum intensity and the Stokes images (top). The maps are produced by averaging a small wavelength region in the continuum for intensity map and in the blue wing of the Stokes profiles for Q, U, and V-maps. The bottom image shows the Stokes spectrum for the slit represented as a dark vertical line in the continuum intensity image at the left.

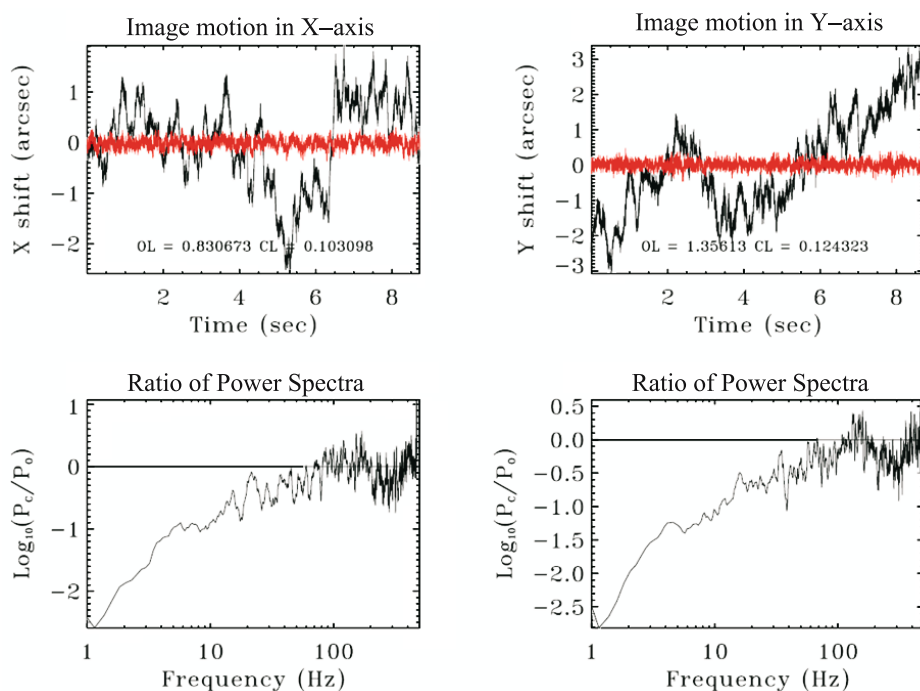


Figure 4. A closed-loop tip-tilt correction of the USO-AO system. A closed-loop update rate of 1 kHz is used. Achieved closed-loop bandwidth is about 100 Hz and the improvement in the 'rms' image motion is anywhere between 5 and 25.

system. A 15 cm telescope is used as a light feed for the laboratory optical set-up. A portion of the image from this 15 cm telescope is selected, by using a field stop, for imaging on to a science as well as a wave-front sensing camera after going through a tip-tilt mirror and a plane mirror (later this plane mirror will be replaced with a deformable mirror). The wave-front sensing camera can operate at a rate of 1 kHz. The algorithm for estimating the image motion is implemented using MMX instructions and a PID type control system is developed for the tip-tilt mirror control. The system was tested in the laboratory using a laser source as well as using the solar image. Closed loop tip-tilt correction was achieved and Fig. 4 shows the performance of the system during a closed-loop operation. Experiments are going on to include the deformable mirror and a closed loop AO correction with the DM is expected to happen in the middle of 2007.

4. Future needs

The overwhelming scientific observations using the AO effected several solar observatories (IRSOL, THEMIS, GREGOR, NST & USO) to develop an AO system for their respective telescopes and they are expected to be functional over the next few years. The two newly proposed large telescope NLST (this proceedings) and ATST (Wagner *et al.* 2006) will require a system with an order more compared to the one currently operating. ATST has proposed for an AO system with 1280 sub-apertures and 1369

actuators and expected to be operated in a closed-loop bandwidth of 250 Hz (Rimmele *et al.* 2006b).

One of the major limitation of the currently operating AO systems is the inability to achieve diffraction limited imaging over a larger field-of-view (FOV). The image becomes blurred away from the isoplanatic patch and the size of this patch (usually about 10") can vary from site to site and also from day to day. The difficulty in consistently locking on to the low-contrast features like granulation is the second major disadvantage with the current AO systems. The faster evolution of the granules compared to high-contrast structures poses this difficulty. The performance of the AO reduces when the observations are done close to the limb and the current AO cannot be used for coronal observations.

Multi-conjugate Adaptive Optics (MCAO) has been proposed as a technique to increase the FOV over which the corrections are done. The Sun is an ideal object to carry out the MCAO due to its extended nature. Solar MCAO efforts are currently underway at the NSO and at Kippenheuer Institute of Sonnenphysik (KIS). There are considerable progress made during the last two years in realising the MCAO (Rimmele *et al.* 2006c; Berkefeld *et al.* 2006).

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SphinX: A Fast Solar Photometer in X-rays

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Abstract. The scientific goals and construction details of a new design, Polish X-ray spectrophotometer are given. It will be incorporated within the Russian TESIS X and EUV complex aboard the forthcoming *CORONAS* solar mission. *SphinX* (Solar **P**hotometer **i**n **X**-rays) will use PIN silicon detectors for high time resolution (0.01 s) measurements of the solar spectra of quiet and active corona in the range 0.5–15 keV. A new filter-fluorescence target concept will be employed to allow for a fast photometry of the solar X-ray flux variations in selected, well defined narrow spectral bands including the Fe XXVI and Fe XXV iron line groups.

Key words. Sun: X-rays—spectrophotometer—high time resolution.

Measurements of the X-ray variability of the Sun are of basic importance for studies of the activity, flares and the space weather. The bulk of coronal thermal plasma of temperatures between 1 MK and 50 MK contributes to the emission in the soft X-ray band. Observations of the spectral variability in the range from 0.5 keV to 15 keV provide the basis for determinations of coronal average temperature, emission measure and related thermodynamic characteristics for non-active corona and flares in particular. For more than 30 years, *GOES* X-ray ion chambers observed the Sun in two spectral bands 0.5–4 Å and 1–8 Å (3–24 keV and 1.5–12 keV respectively). These continuous measurements constitute the reference database as concerns recent solar activity observations. However, they lack the information at both the low- and high-activity conditions (below A and above X10 class). The other active solar spectrometer presently in orbit (*RHESSI*) lacks sensitivity at the lower energy end, below 3 keV. At present low activity conditions, the X-ray signal measured by *GOES* and *RHESSI* does not show-up often above the background. It is of great importance to have a possibility of taking precise measurements of the solar variability in the soft X-ray range, also during periods of very low activity. Therefore, the Wrocław Solar Physics Division Team at the Space Research Centre developed a concept of sensitive spectrophotometer **SphinX**, similar to the Indian SOXS experiment (Jain *et al.* 2005). *SphinX* has recently been incorporated within the TESIS complex, a part of the *CORONAS* payload. The scientific objectives of this instrument are the following:

- Investigations of quiet corona heating processes via photon arrival time–distance analysis.
- Studies of soft X-ray oscillations in the period between 1 s and 500 s.
- Search for transient ionization plasma effects which may provide a tool for determination of flaring plasma densities.
- Determination of coronal plasma composition and its variability for the following elements: Mg, Al, Si, S, Ar, Ca and Fe.
- Determination of differential emission measure (DEM) and studies of its time variability.
- Establishment of soft X-ray reference photometric standard with 5% absolute and 1% relative accuracy.
- Proof of the novel narrow-band photometry (NBF) concept.

SphinX incorporates four PIN, pure silicon, 500 μm thick detectors (AMPTEK XR-100CR) equipped with Be 12.5 μm filters. Three of them with apertures of 19.6, 0.4 and 0.078 mm^2 measure the solar X-ray fluence from events of the *GOES* class as small as 100 times below A1 level, through typical very quiet conditions, i.e., A1 level (rate: 400 cts/s in the largest aperture) up to X40 (rate: 10^5 cts/s in the smallest aperture) over 256 energy bins. The dynamic range thus covers 7 orders of magnitude with the statistical accuracy better than 1% for activity conditions above *GOES* A5.0 level (c.f. Fig. 3). Individual photon arrival times are to be measured to within 1 μs and the spectra are to be taken up to 100 times per second where statistics allow. The detectors are mounted on a thermally isolated support linked through appropriate heat pipe to the external satellite heat radiator. The detectors are by themselves Peltier-cooled. The detectors' operational temperature is expected to be below -50°C . At these temperatures, the energy resolution of the PIN detector is well below 200 eV, allowing for the solar line groups of Mg, Al, Si, S, Ar, Ca and Fe to be distinguished above the continuum emission. The in-flight calibration of the detectors gain is achieved by a special system of the input aperture shutter (see **d** and **e** in Fig. 1). This system, by the command from the SphinX computer moves the shutter sideways, closing detectors' direct solar illumination channels and exposing them to *indirect* fluorescence illumination coming from the back side of thin input transmission filters made of different materials. These three materials are fluorescing at three, precisely known, well separated in energy, monochromatic wavelengths illuminating at the same time the three detectors. The amplitude positions of the individual peaks are being determined and the on-board computer makes appropriate gain corrections if necessary. The widths of the individual peaks contain information on the energy resolution of the detectors. Depending on the intensity of the solar illumination, a secondary, pile-up peaks can be resolved, creating thus a possibility to dynamically calibrate also this effect. The calibration sequences can be run any time, and depending on the level of the solar illumination causing the input filter fluorescence, the duration of the calibration will take any time down to few seconds. After the calibration, the shutter is moved to the original position, where a direct illumination of the detectors through their respective apertures is resumed. In such gain calibration solution, the detectors have always three energy-amplitude reference points available for calibration (to the accuracy of 5–7 eV) and no radioactive sources are to be used.

The fourth detector, also connected to the heat pipe, measures instantaneous narrow-band fluxes in three wavelength bands: 1.74–2.07 $\text{\AA}$, 2.50–3.08 $\text{\AA}$ and 7.95–9.54 $\text{\AA}$

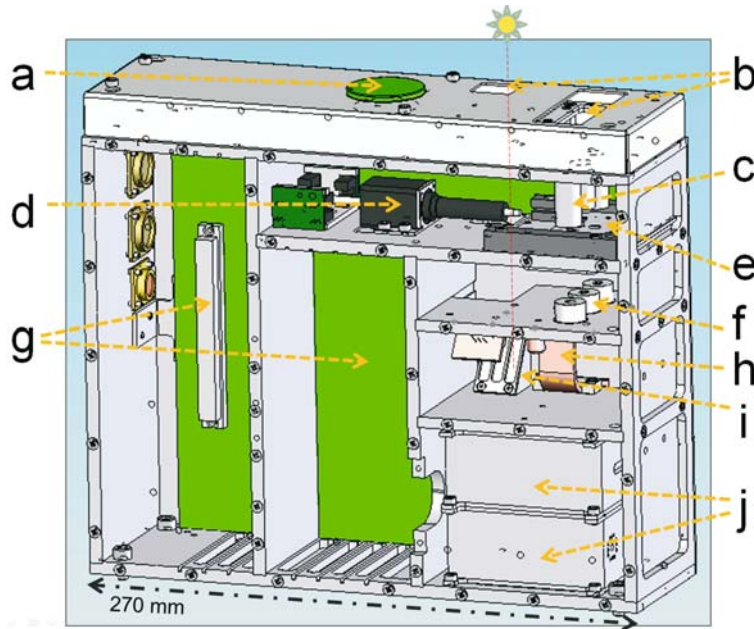


Figure 1. The design of SphinX with main components indicated: (a) alignment mirror, perpendicular to the Sun direction, (b) entrance openings covered with optical EUV stoppers – example solar ray is indicated, (c) mechanical collimators narrowing the detectors FOV to within $\pm 2.5^\circ$ of the Sun, (d) the shutter motor, (e) moving tray with the shutter and fluorescence calibration filters, (f) three spectrophotometric detectors, (g) electronics with microcontroller (h) cooling heat pipe, (i) three narrow band filter-detector sections and (j) pre-amplifiers, independent for each detector. The horizontal extension is indicated at the bottom.

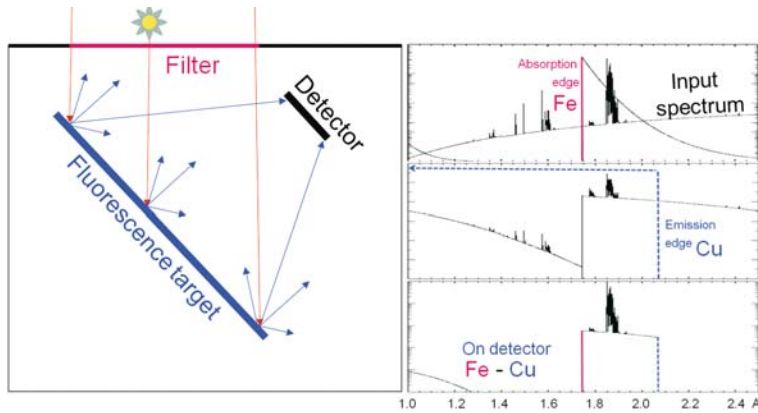


Figure 2. (Left): The concept of narrow band filter (NBF); (right): solar X-ray spectrum transmitted through the filter contains predominantly the radiation above the wavelength of specific absorption edge (middle panel). This illuminates the fluorescence target and only those wavelengths below the fluorescence excitation edge contribute to the scattered flux measured by the detector. In respect with, the system acts as the narrow-band filter with precisely defined, naturally sharp energy boundaries.

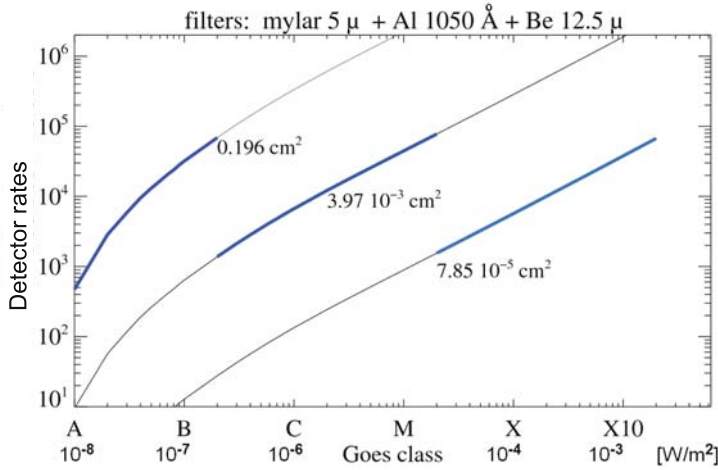


Figure 3. The expected count rates for SphinX detectors for the three apertures to be used. It is to be ~ 400 cts/s at the low solar activity A1 (10^{-8} W m^{-2}), through the largest aperture. The expected orbital background rate is ~ 10 cts/s. The ranges of useful performance for each of the detectors are denoted by thicker portions on the curves. Even for events of GOES class X40 ($4 \cdot 10^{-3} \text{ W m}^{-2}$), the signal will not be saturated on the smallest aperture detector.

up to hundred times per second using novel measurements concept presented in the right panel of Fig. 2. In this NBF concept, three fluorescing targets illuminate a single PIN detector. The fluorescence radiation is coming from three different pure element targets being illuminated through separate, individual selected input filters made of elements of slightly higher atomic number than their respective targets. Therefore, on the detector, the three emission peaks of different amplitudes are being instantaneously recorded. Their relative strengths depend on the shape of the solar incident spectrum in three narrow spectral bands and therefore are very sensitive to plasma temperature. We have selected the filter-target materials in order to record the portions of solar spectra containing strong emission lines of iron and magnesium. This will allow us to study transient effects, expected to be seen when emitting plasma conditions vary with time scales less than few seconds. The overall absolute calibration of the detectors has been performed using Berlin BESSY (www.bessy.de) synchrotron beams. The absolute calibration accuracy against the synchrotron beam is expected to be better than 5%. The synchrotron source has a spectrum similar to the solar one so we envisage SphinX to be the best calibrated soft X-ray spectrometer ever flown. In addition SphinX has been calibrated at several monochromatic wavelengths using the Palermo XACT facility (Barbera *et al.* 2006). The general SphinX characteristics are as follows: dimension – $270 \times 70 \times 220$ mm, weight – 3.5 kg, power consumption – up to 15 W, 5 W average, overall data rate – more than 20 MB/24^h, limited only by the on-board computer storage and telemetry resources. The instrument is equipped with its own microcomputer, reading out in-flight performance (detector temperatures, motor and shutter positions, voltages, etc.) and issuing various flare flags to its “mother” TESIS computer and to the other instrument aboard CORONAS satellite. Extended SphinX instrument paper is to be published in *Solar Physics* shortly.

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Proposed National Large Solar Telescope

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Abstract. Sun's atmosphere is an ideal place to study and test many magnetohydrodynamic (MHD) processes controlling turbulent plasma. We wish to resolve some of the finest solar features (which remain unresolved presently) and study their dynamics. Indian Institute of Astrophysics has proposed to design, fabricate and install a 2-meter class solar telescope at a suitable site in India to resolve features on the Sun of the size of about 0.1 arcsec. The focal plane instruments will include a high resolution polarimetric package to measure polarization with an accuracy of 0.01 per cent; a high spectral resolution spectrograph to obtain spectra in 5 widely separated absorption lines simultaneously and high spatial resolution narrow band imagers in various lines. The Himalayan region appears to be a good choice keeping in view the prevailing dry and clear weather conditions. We have started detailed analysis of the weather conditions in the area and at some other locations in India. The site characterization will be done using the Sun-photometer, S-DIMM and SHABAR techniques to determine the seeing conditions.

Key words. Solar telescope—adaptive optics—spectropolarimetry—location.

1. Introduction

The Sun offers an excellent laboratory to address the key processes that take place in highly magnetized plasma that govern the many astrophysical phenomena and thus provides a stepping stone to infer physical parameters in other more complex stellar systems and of the universe at large. All the observatories in India (Kodaikanal, Nainital and Udaipur) have made significant contribution to solar physics. Kodaikanal observatory has obtained photoheliograms and Ca-K and H- α spectroheliograms of the Sun over a period of about 100 years and formed unique datasets to study the long period variations in various physical phenomena on the Sun because of uniformity in the data (Singh & Bappu 1981; Singh & Prabhu 1985; Sivaraman *et al.* 1999; Makarov *et al.* 2001). Apart from the study of long period variations, one needs to investigate very small scale structures such as flux tubes that play an important role in all the activity and physical processes on the Sun. One meter Swedish vacuum solar telescope at La Palma (a beautiful sight with very stable and transparent sky conditions), Canary Islands which is being operated for the last 5 years by the Royal Academy of Sciences,

Sweden has yielded great details about the solar surface features (Scharmer *et al.* 2002) of the order of 0.3 arcsec. The 65-cm telescope at Big Bear Solar Observatory has been used to take images and make polarization measurements in 1565 nm line recently (Cao *et al.* 2006a, 2006b). They could achieve a spatial resolution of 0.3 arcsec using adaptive optics. Kiepenheuer Institute of Solar Physics, Germany is planning Gregorian, on-axis, Alt-Azimuth, 1.5 m open telescope to be installed at Izana. Big Bear Solar Observatory is planning to fabricate and install 1.6 m off-axis in place of 65 cm existing telescope. The upgraded facility will be used for the existing programs and magnetic field measurements with better spatial resolution and photometric accuracy.

A large solar telescope known as Advanced Technology Solar Telescope (ATST) of 4-meter size, capable of taking diffraction limited images of the Sun (spatial resolution of 0.03 arcsec) and make observations in the infrared wavelength region, has been proposed by the National Solar Observatory in collaboration with many organizations (ATST report 2000). The proposed telescope will be capable of observing the solar corona also. It is likely to come up in the near future. There are many scientific objectives of the telescope:

- How are the highly intermittent magnetic fields observed at the solar surface generated by dynamo processes and how are they dissipated?
- What magnetic configurations and evolutionary paths lead to flares and coronal mass ejections?
- What mechanisms are responsible for variations in the spectral and total irradiance of the Sun and solar-type stars?
- Progress in answering these critical questions requires a study of the interaction of the magnetic field and convection with a resolution sufficient to observe scales fundamental to these processes. Recently a 50-cm telescope known as Hinode has been launched in space by a Japanese group in collaboration with USA and UK scientists and has obtained diffraction limited images with a spatial resolution of 0.2 arcsec in G-band, Ca-H line and have made magnetic field measurements (Hinode 2006). The images indicate that flux tubes still need to be resolved by making observations with higher spatial resolution. ATST is planning to achieve this and has many other objectives that require a large size aperture of about 4 m. We, in India plan to have limited scientific objectives such as to resolve the flux tubes, make magnetic field measurements with required accuracy and some other objectives listed in Table 2 for which a 2 m class telescope seems to be sufficient.

It is well known that existing observational facilities for solar research in India, the latest being Solar Tower Telescope installed in the year of 1960, are grossly inadequate for high resolution observations of the Sun. Absence of a large and versatile telescope, which can facilitate simultaneous measurements of the solar atmospheric parameters and of the vector magnetic fields with high accuracy, has been a serious handicap for the solar astronomers. Efforts have been made for the past 20 years or so to have a good and large solar telescope, equipped with state-of-the-art focal plane instruments using the present day technology but without success. Recently Udaipur Solar Observatory has succeeded in designing a moderate size telescope and placed an order with a company (Venkatakrishnan 2006). There is therefore, an imperative need for a state-of-the-art solar observational facility, comparable to the best in other countries. A 2-meter class

large solar telescope is hence being proposed as a national facility for implementation during the 11th five year plan.

2. Scientific objectives

The proposed telescope will address the fundamental questions about the nature of solar magnetism, will aim to resolve flux tubes and measure their strength; address the development of magnetic fields on the Sun which are responsible for almost all the observable phenomena on the Sun such as solar dynamo, solar cycle and solar variability that determine and control the space weather. The other scientific objectives are: (i) MHD waves by resolving small structures and determining periods of oscillations which may be responsible to transport the energy to the upper atmosphere of the Sun; (ii) Dynamic evolution of small scale structures by making high cadence observations; (iii) Evolution of active regions and their role in triggering solar flare, prominences, filament eruptions, CMEs, etc.; (iv) Thermodynamics of the chromosphere by making the observations in the infrared wavelengths; (v) Weak and turbulent magnetic field measurements using Hanle effect which are as important as the strong magnetic fields and have now become possible to be measured because of the development in technology. All these objectives will be achieved by making observations with high spatial resolution using adaptive optics, high spectral resolution, high temporal resolutions, multi-wave length capability of imaging and spectroscopy focal plane instruments, high photon flux and sensitivity of the detectors and using the infrared part of the spectrum for observations. Table 1 gives a summary of the scientific objectives.

3. Telescope specifications

Keeping in view the scientific objectives and technical requirements of observations, it has been proposed to have a 2 m class telescope of Gregorian format. Direct light from the Sun falls on the primary mirror M1 of aperture 2 m and of a focal ratio of $f/2.0$ which forms a solar image of about 36 mm size at the prime focus F1. In the process of forming the solar image, the primary mirror concentrates nearly 3 kW of heat at the focus. A heat trap (not shown in Fig. 1) with an efficient circulating coolant placed close to F1 takes away more than 98% of the heat by directing most of the solar radiations out of the telescope to the atmosphere. A field stop with a hole of about 6 mm diameter allows only a small portion of the solar image (about 300 arcsec) to pass through further to the secondary mirror M2. This helps to reduce the heat further within the telescope. The secondary mirror M2 magnifies the 300 arcsec of the Sun's image and brings it to the secondary focus F2. The beam is allowed to proceed further to the tertiary mirror M3 of the Gregorian system which magnifies the image at the secondary focus and brings the 300 arcsec image of the Sun to the final focus point F3 via a chain of flat mirrors M4, M5, M6, M7, and M8 into the laboratory at the lower level as shown in Fig. 1. M6, M7, and M8 form a de-rotator that compensates for the rotation of the field of view. Mirror M9, serves as the tip tilt mirror (active optics mirror) and M10 is the deformable mirror DM of the adaptive optics. The mirror M11 feeds light to the chosen focal plane instruments for various observational programs through set of beam splitters and mirrors. The primary image formed by the mirror M1 will have aberrations (coma and astigmatism). These are

Table 1. Summary of science objectives and matching instrumentation.

Summary of science objectives				
Physical process/ region to be observed	Physical quantity to be measured	Telescope/instrument requirement	Instruments proposed	
MHD waves and oscillations	Intensity variation of 1% or less	High photon flux needing aperture of 2 meter and above	Narrow band filters	
	Velocities of ~ 200 m/sec or less	Spectral resolution of a few mÅ	High resolution spectrograph	
	Magnetic field measurements	Polarization accuracy: 10^{-4} or better	Spectropolarimeter with spatial information	
	Properties of oscillations	High time cadence	High speed cameras	
Different heights in solar atmosphere	Diagnostic parameters covering many spectral lines	Both visible and IR capabilities	Ca II K, H α , G band, He I 1083 nm, Fe I 1.56 microns filters	
Active region evolution	Velocity and magnetic fields	High time cadence large field of view of 300 arcsec	Spectropolarimeter with spatial information	
Hanle effect	Magnetic field ~ 5 –10 G	Polarization accuracy 10^{-4}	Spectropolarimeter	

properly corrected by mirrors M2 and M3 to produce the final aberration-free image of the 300 arcsec of the Sun at the final focus. In addition, these two mirrors will be mounted on hexapods to enable fine tuning of the alignment of the telescope system. The f-ratio of the telescope system will be designed to be f/40 so that the final magnified image of the Sun will have a scale of ~ 2.5 arcsec/mm at the final focal plane.

One of the major focal plane instruments will be a spectropolarimeter package to measure the magnetic fields with a high degree of accuracy and high spatial resolution. The spectropolarimeter will use the new technology of Dense Wave Division Multiplexing (DWDM) filters and Fiber bundles to convert two dimensional solar images into slit images to perform spectroscopy. This will avoid scanning of the image to measure the magnetic fields and thus permit study of the temporal variations in the magnetic fields before, during and after the solar flares. This will help us to understand the processes involved in the triggering of flares. A high spectral resolution spectrograph capable of taking spectra in 5 widely separated absorption lines simultaneously and a facility to take high spatial resolution narrow band images of the Sun in many photospheric and chromospheric lines to resolve structures of about 50 km size. The optical layout of the telescope is shown in Fig. 1 and the specifications are indicated in Table 2. The narrow band filters will include Ca-K, H- α , CN-band, G-band and UBF with 0.02 nm pass band. The telescope will employ active and adaptive

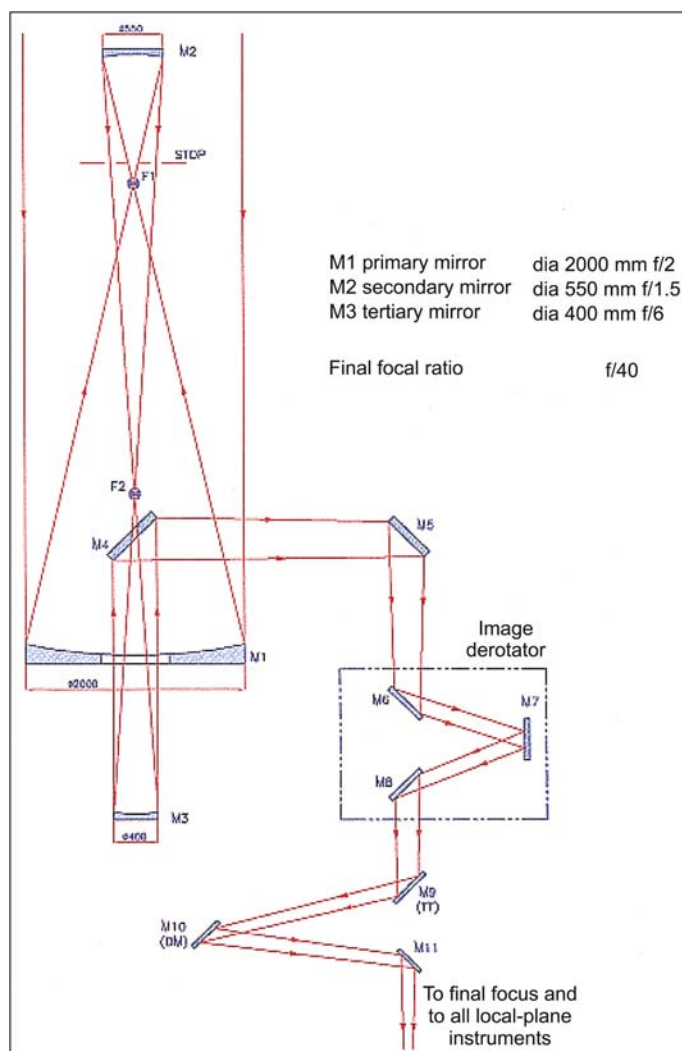


Figure 1. Optical layout of the proposed 2 m solar telescope.

optics to stabilize the image and reduce the effect of variations in the atmospheric conditions.

4. Location and site survey

The location of the telescope needs to provide a large number of clear hours for making observations with very good seeing and transparency. To make observations in the infrared wavelengths for high accuracy of magnetic field and velocity measurements, water vapours in the air need to be very less. Himalayan regions appear to provide such atmospheric conditions. A 2 m telescope has already been established at Hanle to perform astronomical observations during night-time. The site survey done at various

Table 2. Summary of telescope features.

Aperture (primary mirror M1)	2 meters
Focal length	4 meters
Optical configuration	3 mirror, Gregorian on - axis
Field of view	300 arcsec
Final focal ratio of the system	f/40
Image scale	2.58 arcsec mm ⁻¹
Optical quality	< 0.1 arcsec over the field of view
Wavelength of operation	380 nm to 1000 nm (upto 2500)
Polarization accuracy	1 part in 10,000
Active and adaptive optics	To realize near diffraction limited performance
Spatial resolution	< 0.1 arcsec

**Figure 2.** Part of the Pangong lake in the Himalayan region.

places has indicated that the lake sites provide better seeing conditions such as that at Big Bear Solar Observatory. We, therefore, have planned to determine the atmospheric conditions during daytime at Hanle (existing observatory), Pangong in the Himalayan region (a big lake site, Fig. 2) and Devasthal near Nainital Observatory). We have already collected the weather data at Hanle and at Devasthal. We have started to determine the weather conditions at Pangong as shown in Fig. 3. We plan to measure the seeing conditions at these places using S-DIMM and SHABAR techniques. We have already procured one such instrument from National Solar Observatory, USA which we expect to install at Hanle during the period of June 2007. We plan to fabricate such systems to measure the seeing conditions at other places mentioned above and at some other places chosen later.



Figure 3. Weather station at Pangong to measure solar radiations, temperature, wind speed, etc.

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Development of a Low-order Adaptive Optics System at Udaipur Solar Observatory

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Abstract. A low-order Adaptive Optics (AO) system is being developed at the Udaipur Solar Observatory and we present in this paper the status of the project, which includes the image stabilization system and calibration of wavefront sensor and deformable mirror. The image stabilization system comprises of a piezo driven tip-tilt mirror, a high speed camera (955 fps), a frame grabber system for sensing the overall tilt and a Linux based Intel Pentium 4 control computer with Red Hat Linux OS. The system operates under PID control. In the closed loop, an rms image motion of 0.1–0.2 arcsec was observed with the improvement factor varying from 10–20 depending on the external conditions. Error rejection bandwidth of the system at 0 dB is 80–100 Hz. In addition to that, we report the on-going efforts in the calibration of lenslet array and deformable mirror for sensing and correcting the local tilt of the wavefront.

Key words. Sun: astronomical seeing, adaptive optics, solar telescopes, solar imaging.

1. Introduction

A Multi-Application Solar Telescope (MAST) equipped with an Adaptive Optics (AO) system is proposed to be installed at the lake site of Udaipur Solar Observatory (USO) in India. It is planned to build the AO system for high resolution observations in two phases. The first phase involves the development of an image stabilization system (Sridharan *et al.* 2005), which will correct the global tilt of the distorted wave-front and deliver a steady image. The second phase involves correction of local tilts of the residual wavefront using a deformable mirror.

The image stabilization system is in operation and provides an improvement of 10–20 times in the rms image motion. Now, the second phase is in progress. In this paper, we present some of the results we achieved with our image stabilization system and the ongoing efforts for achieving local tilt correction of the wavefront.

2. Image stabilization system

2.1 Image stabilization system: Global tilt correction

In an image stabilization system, a small portion (typically 10×10 arcsec²) of a short exposure (typically < 10 ms) image is cross-correlated with a previously recorded reference image. Subsequent images are correlated with the same reference image. The reference image is updated typically every 10 seconds by the best image selected during that period. The cross-correlation extracts the shift of the instantaneous image with respect to the reference image in two orthogonal directions. These shifts are used to feed appropriate voltages to the tip-tilt mirror so that its surface closely approximates the instantaneous global tilt in the wave-front. As the global tilt is removed from the wave-front after being reflected by the tip-tilt mirror, the image motion is arrested. This results in a stable output image in the science camera and thus allows for long exposures.

2.2 Description of optical set-up of image stabilization system

The image formed by the 15 cm Coude refractor (f/15 beam) is re-imaged by a combination of two doublets inside the laboratory with a magnification of two. A portion of the image is collimated by a doublet of focal length 200 mm. The collimated beam is reflected by the tip-tilt mirror and another plane mirror and is finally focused by a lens of focal length 300 mm. A beam splitter splits the converging beam equally into two parts: one beam is directed to a ‘science camera’ after passing through a filter of 1 nm bandwidth centered at 430.5 nm and the other beam is directed to the ‘wave-front sensing camera’ after passing through suitable neutral density filters. The final image has a scale of 0.31 arcsec per pixel at the ‘sensing camera’. The light feed and schematic of the optical set-up are shown in Fig. 1.

2.3 System components

Apart from optical elements, the main components of the system are: (1) A 16-bit 2-channel digital-to-analog converter board, (2) the piezo-electric stage and its controller, and (3) the data acquisition system which consists of a fast camera (955 fps) and a frame-grabber card. These are similar to those used in the low cost AO system at NSO/Kitt Peak, USA (Keller *et al.* 2003, hereafter KPA).

2.4 Software

The AO system software consists of two C-programs that are executed together and share a common memory through inter process communication mechanism. The first one called “the universal tracker” is largely based on the Linux device driver of the frame grabber card provided by *GOM mbH*. The second one called “the control software”, is used to control the various parameters of the system, including PID. At USO, we have modified the aforesaid software developed by KPA to suit our system. The image stabilization system at USO is working with the closed loop update rate ≈ 1 kHz. The closed loop correction bandwidth is ~ 80 – 100 Hz as shown in Fig. 2. The rms image motion in the closed loop is about 0.1 arcsec and is better than that in the open loop operation by a factor of 10–20.

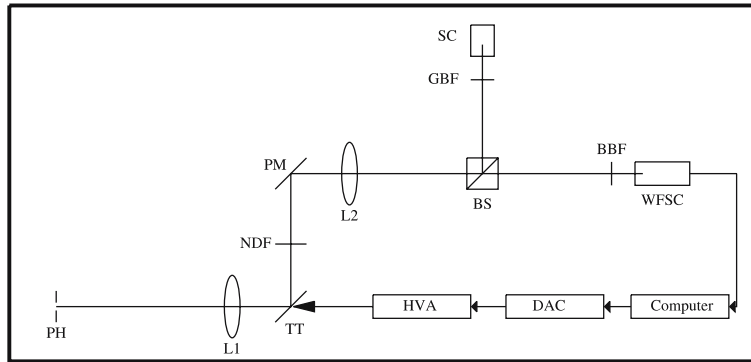
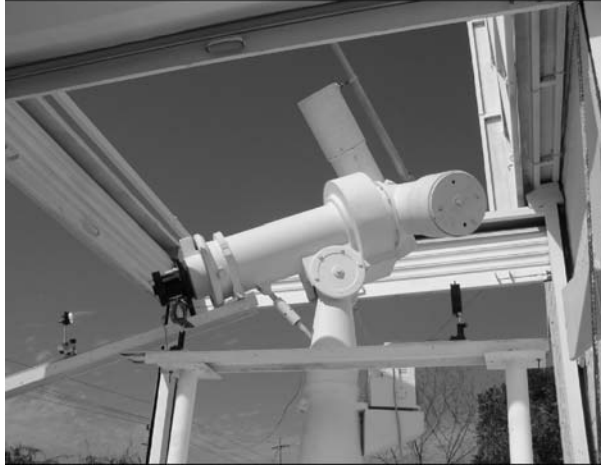


Figure 1. Coude telescope (top panel) at the light feed and schematic of the optical set-up (bottom panel) inside the laboratory is shown here. PH: Field stop, L1 and L2: Lenses, PM: Plane mirror, BS: Beam splitter, WFSC: Wavefront sensing camera, SC: Science camera, PC: Control computer, HVA: High voltage amplifier, DAC: Digital to analog converter.

3. Correction of local tilt of the wavefront

Correction of the local tilt of the wavefront involves two important steps, the first one being sensing the local tilts and second one is correction. We employ Shack-Hartmann Wavefront Sensing (SHWFS) using a lenslet array (9×9 lenslets or sub-apertures) and a 37-channel micro-machined membrane mirror (DM) for correction. The pitch and the focal length of the lenslet array are $600 \mu\text{m}$ and 145 mm , respectively. This pitch size of lenslet array and CCD allows us to accommodate only 17 sub-apertures in the sensing camera. The code for calculating the local tilts of the wavefront using the lenslet array has been successfully tested (Bayanna *et al.* 2005). The DM and lenslet array were characterized through lab simulations before being used for solar observations. An optical set-up was made using a He-Ne laser source inside the lab. Using suitable lenses and apertures, we obtained a 15 mm collimated beam from the laser source. Two similar DMs were employed in the set-up to use one of them as a source of distortion and the other as a correcting element. However, in this paper

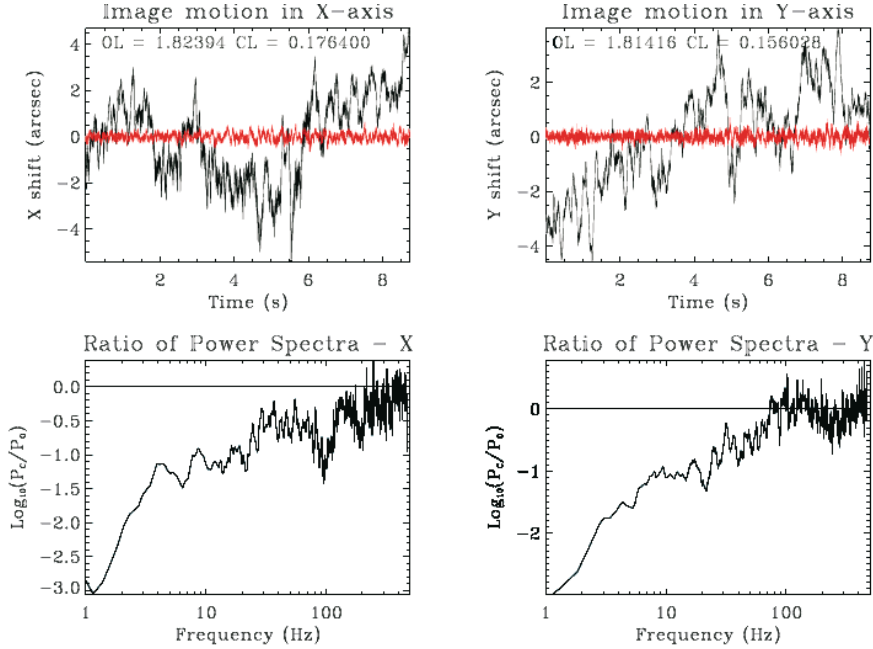


Figure 2. Top panels show image motion in the open loop (OL) and closed loop (CL) in black and red, respectively. Bottom panels show the ratio of power spectra of closed to open loop. P and D control parameters have been used in tuning the system. The values for *Channel 0* are: $P_0 \approx 20$ and $D_0 \approx 80$ and that for *Channel 1* are: $P_1 \approx 0.45P_0$ and $D_1 \approx 0.45D_0$.

we are presenting the calibration details only, which has been completed. Before calibrating the DM, mapping of the actuator space with the sub-aperture space was carried out, so that the distortion could be created at the desired location and in a systematic manner. The main purpose of the calibration was to estimate the influence matrix of the deformable mirror, i.e., how a change in voltage given to a particular actuator affects the change in shift in all the sub-apertures. Voltage was applied to each actuator of the DM from 0–225 DACs and the corresponding image motion in each sub-aperture was recorded. The slopes of voltage vs. shift curves for all the sub-apertures were estimated. The same procedure was repeated for all the actuators. The influence matrix is a rectangular matrix of size $2M \times N$, with the number of columns equal to the number of actuators (N) and the number of rows equal to the number of slope measurements ($2M$), which is twice the number of sub-apertures. The matrix is inverted using Singular Value Decomposition. As we have 17 sub-apertures and 37 actuators, to have practical and best suited solutions by inverting the influence matrix, we restricted our usable actuators to only 7. To check the influence matrix and the code for the estimation of voltages from the shift vector, an arbitrary voltage set was applied to the deformable mirror and the shift vectors along the x - and y -directions in the sub-aperture plane were recorded as shown by the solid line in Fig. 3. Using the observed shift vector and the influence matrix, a voltage vector was obtained, which was in agreement with the applied voltage vector. The application of this voltage vector produced nearly the same shift vector as the one corresponding to the applied voltage set. These are shown by the

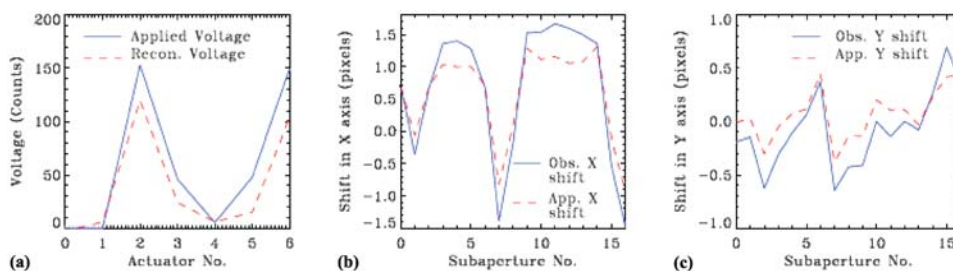


Figure 3. (a) Plot depicts voltages applied to the deformable mirror and the reconstructed voltage set obtained from the influence matrix. (b) and (c) Plots depict the shift obtained for a given voltage set and the shifts obtained from the reconstructed voltage set along x and y axes respectively.

dashed lines in Fig. 3. During this process, we have observed that in the present optical set-up the DM is insensitive at voltages smaller than 60 DACs. This is why the shift observed and the consequent voltage vector obtained through the influence matrix are very small for low values of applied voltages. We also observed that for the image motion more than 0.5 pixels, the voltage vector obtained is sensitive enough to influence the DM surface. Thus, even if the image motion in one of the sub-apertures due to the local wavefront tilt is above 0.5 pixels, the subsequent voltage vector will be sensitive enough to suppress it to the sub-pixel domain, which is our requirement. The DM is ready for integration with the image stabilization system and for dynamic correction for complete realization of the AO system.

Acknowledgements

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CONCEPTUAL PAPER

Utilization of GPS Satellites for Precise Irradiation Measurement and Monitoring

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Abstract. Precise measurement of irradiance over the earth under various circumstances like solar flares, coronal mass ejections, over an 11-year solar cycle, etc. leads to better understanding of Sun–earth relationship. To continuously monitor the irradiance over earth-space regions several satellites at several positions are required. For that continuous and multiple satellite monitoring we can use GPS (Global Positioning System) satellites (like GLONASS, GALILEO, future satellites) installed with irradiance measuring and monitoring instruments. GPS satellite system consists of 24 constellations of satellites. Therefore usage of all the satellites leads to 24 measurements of irradiance at the top of the atmosphere (or 12 measurements of those satellites which are pointing towards the Sun) at an instant. Therefore in one day, numerous irradiance observations can be obtained for the whole globe, which will be very helpful for several applications like Albedo calculation, Earth Radiation Budget calculation, monitoring of near earth-space atmosphere, etc. Moreover, measuring irradiance both in ground (using ground instruments) and in space at the same instant of time over a same place, leads to numerous advantages. That is, for a single position we obtain irradiance at the top of the atmosphere, irradiance at ground and the difference in irradiance from over top of the atmosphere to the ground. Measurement of irradiance over the atmosphere and in ground at a precise location gives more fine details about the solar irradiance influence over the earth, path loss and interaction of irradiance with the atmosphere.

Key words. Albedo—global positioning system—solar irradiation.

1. Introduction

Solar irradiance is the driving factor behind the earth's life cycle. Any small change in solar irradiance will have a straight forward effect on the earth. Measuring and monitoring the precise irradiation will lead to a better understanding of the Sun–earth relationship. The irradiation variation over earth surface and on top of the atmosphere is very small. But that very small change will play a crucial role in the earth's radiation budget and also affect the earth climate. The well-known reason for the irradiation

variation is that the Sun is not a constant source of radiator. The Sun keeps on changing throughout its lifetime and these variations are due to solar flares, coronal mass ejections, solar cycle variations, etc. Until now the irradiance measurements are made at several places and these are inter or extrapolated to obtain the global irradiance over earth and on the surface. To obtain the precise irradiation value over the earth irradiation is to be monitored over the top of the atmosphere and at the earth's surface simultaneously. The reason to monitor the irradiation over top of atmosphere and on the surface of the earth at the same point is to enable monitoring the irradiation loss during the atmosphere transmission and also to understand small changes in the irradiation.

2. GPS

Global Positioning System is a constellation of 24 satellites which orbit at an altitude of 20,200 km above the earth's surface in an inclined elliptical orbit. GPS provides very precise latitude and longitude details which will be in centimetres. GPS covers the whole earth and any point on the earth surface can be pin-pointed. The reason for choosing GPS satellite is because of its continuous availability throughout the year. If the future GPS satellites like GLONASS, GALILEO, etc., are equipped with irradiation measuring instruments like Total Irradiance Monitor [SORCE], etc., and moreover, if the measurements are made continuously over the earth's atmosphere then the amount of irradiation data available are enormous which will lead to better results in irradiance monitoring. During any particular year at least a minimum of 12 satellites will be facing the Sun and monitoring the irradiance over that region and the orbital paths of the satellite will be well-known. One more added advantage of using the GPS satellite is the repetitive coverage over an area on a single day, so the variation in the same area over a short period of time can also be calculated.

3. Space- and ground-based measurement and monitoring

At any instant nearly 12 satellites will be facing the Sun and they will measure the irradiance. The orbital path of the satellites are predefined and well-known. To measure the irradiation precisely the measurement should be made at ground also at the same instant of time when the satellite is passing over a certain area. The ground measurements are performed to reveal the solar irradiation available on the ground after several atmospheric path losses. The reasons to measure the irradiation both on top of the atmosphere and on the ground are to understand the changes occurring in the irradiation during their travel to the surface, etc. Ground instruments are required at several clearly available places for measurement.

For example, if the satellite is at a specific location in space and their co-ordinates in terms of latitude and longitude are known by pre-calculations, e.g., $13^{\circ}00' \text{ N } 80^{\circ}00' \text{ E}$, the satellite measures the irradiation at that particular position in space and if the measurements are made in the ground at the same place then we can compare the changes in the values. If more number of ground stations are established then for each ground station data we will obtain the top of the atmosphere data from the satellite. More stations will lead to better results, because we can compare more ground data with the satellite data.

For a particular instant of time and at a particular point we will get the irradiation at the top of the atmosphere, at the ground, the difference in irradiation, the amount of loss

during that time and the variation in irradiation when compared to the corresponding previous readings at the same point of observation. Though the irradiation variation on the top of the atmosphere is very small, monitoring those small variations will also help us to understand the influence of solar radiation on earth. The ground monitoring is difficult also because of the multiple scattering of radiation from the ground which makes the reading difficult. Though there are some difficulties, the data obtained from these calculations will be better than the interpolated or extrapolated values which are widely used.

4. Possible outputs and advantages

- For a single day, numerous data would be obtained from single locations in space and on earth. Averaging these irradiances will lead to accurate irradiance values for a day which would be a better output than either by interpolation or extrapolation techniques or current estimation techniques.
- If the irradiances are monitored for an 11-year solar cycle, then the effects of the Sun on earth can be studied with the real data.
- For precise irradiance calculation and to monitor their changes, the ground and space data should be incorporated which will yield better output of irradiance values.
- Continuous measurement decreases uncertainties and removes random errors.

5. Conclusion

- GPS satellites can measure and monitor solar activities and measurements of one full solar cycle (11 years) will provide us enormous details about the Sun–earth relations.
- Combining ground and satellite data will provide us much more details like path loss, atmospheric variation over a period, irradiation variation, etc.
- To completely understand the Sun–earth relationship, monitoring the solar activity and their effects on earth with the use of GPS satellites and simultaneous monitoring at ground will improve our knowledge in their relationships.

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Error Minimization of Polynomial Approximation of Delta T

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Abstract. The difference between Universal time (UT) and Dynamical time (TD), known as Delta T (ΔT) is tabulated for the first day of each year in the Astronomical Almanac. During the last four centuries it is found that there are large differences between its values for two consecutive years. Polynomial approximations have been developed to obtain the values of ΔT for any time of a year for the period AD 1620 to AD 2000 (Meeu 2000) as no dynamical theories describe the variations in ΔT . In this work, a new set of polynomials for ΔT is obtained for the period AD 1620 to AD 2007 that is found to produce better results compared to previous attempts.

Key words. Delta T —polynomial approximation—Dynamical time—Universal time.

1. Introduction

Currently, three different time scales are in use for modern astronomy as well as for civil life (McCarthy 2004). The Universal time (UT), based on the rotation of earth around its own axis, Ephemeris time (ET), based on the revolution of the Earth in its orbit around the Sun and the Atomic time (AT) based on the quantum mechanics of atom. Each of these bases for measure of time has its history of evolution (Nelson *et al.* 2001).

Due to irregularities in the axial rotation of the Earth, the Universal time is not a uniform time (Husfeld & Kronberg 1996). Currently the Earth is slowing down and its rate is not constant, instead it has unpredictable irregularities (Meeus 1998a). Ephemeris time (ET) is not affected by the irregularities of the axial spin of the Earth, as it is evaluated from the revolution of the Earth in its orbit around the Sun in view of the principles of celestial mechanics. In this context it is clear that ET and UT are not perfectly in agreement (Seidelmann 1985). Moreover, from the theory of relativity we know that time is a dynamical quantity and it could vary from point to point in space (co-ordinated time). In 1984, the ET was replaced by Terrestrial Dynamical time (TDT) (Meeus 1998a), which is defined by an atomic clock located at the origin of a co-ordinate system for which this time is being defined. If the origin of this co-ordinate system is chosen at the centre of the Earth, we have Terrestrial Dynamic time (TDT) or if this refers to the centre of mass of the solar system we have Barycentric dynamical time (TDB) (Guinot & Seidelmann 1998). Later TDT was renamed as “Terrestrial

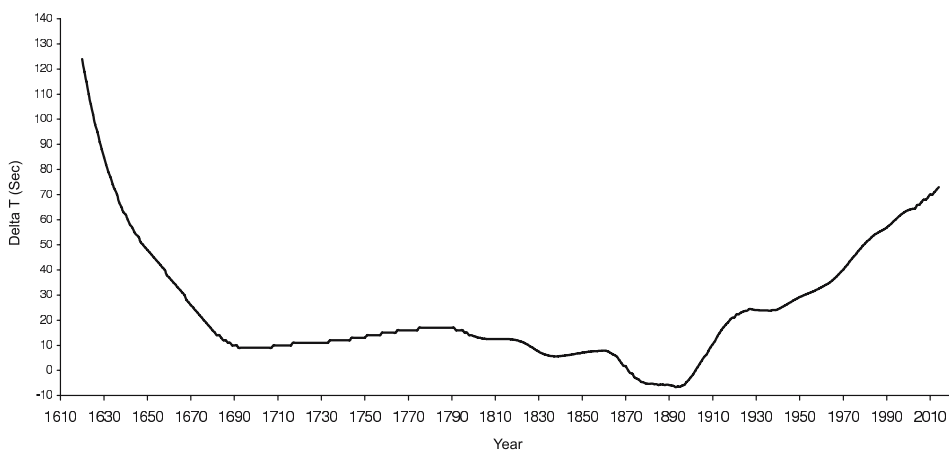


Figure 1. Shows the variation of ΔT with epoch 1620–2007.

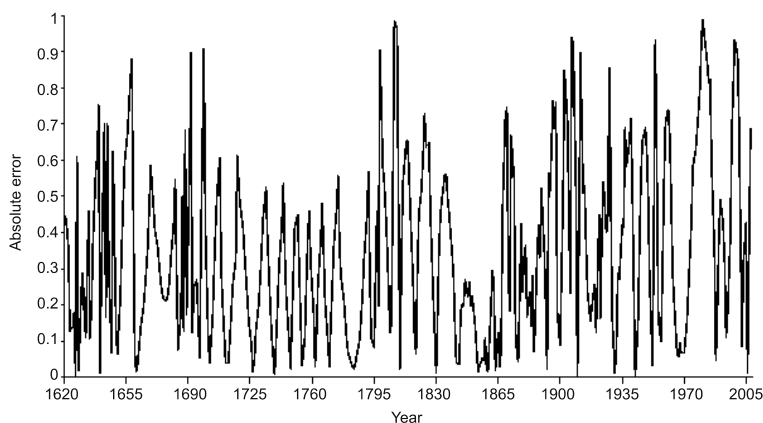


Figure 2. Shows the absolute error of new approximation.

time” (TT) (Nelson *et al.* 2001), which is now considered to be uniform time scale, and is used as the time argument for the prediction of the astronomical events. Our clocks, locked on UT are gradually slowing down with respect to the uniform TT and hence using UT, astronomical events seem to occur earlier than predicted (Seidelmann 1985). The difference between TT and UT (Meeus 1998b) is shown in Fig. 1 (Robert 2002).

2. Polynomial approximation of Delta T by Meeus

For any instant during a year one needs some approximation technique to deduce the value of ΔT . One such technique was developed by Jean Meeus and Larry Simons in 2000 that gives a set of polynomial approximation for ΔT for the range of 1620 to 2000 (Meeus & Simons 2000). It is difficult to find a single polynomial approximation for the whole span of 1620 to 2000. Meeus & Simons have divided the whole data set into eight segments/intervals (Table 1 of Meeus & Simons 2000) and determine a fourth degree polynomial for each segment that approximates ΔT for that segment.

Table 1. Depicts the co-efficient of new polynomial approximation.

Time interval (year)	k	a_0	a_1	a_2	a_3	a_4
1620–1698	3.48	38.067	−105.262	14.523	−273.116	1162.805
1699–1806	2.545	13.759	13.893	7.591	−39.048	−71.724
1807–1872	1.675	5.859	−3.654	161.524	−157.977	1612.55
1873–1906	1.175	−6.203	−2.732	139.921	1006.463	6250.501
1907–1948	0.795	24.006	12.382	−234.449	1055.209	1815.042
1949–2007	0.29	47.917	91.081	−29.979	−358.707	262.919

Using these polynomials one can calculate the value of ΔT at any instant of the year for the range 1620 to 2000 AD. Their technique is based on the formula

$$\Delta T = a_0 + a_1u + a_2u^2 + a_3u^3 + a_4u^4, \quad (1)$$

where ‘ u ’ is given by

$$u = k + \frac{(\text{year} - 2000)}{100}, \quad (2)$$

‘year’ could be a fractional value from 1620 to 2000 and k is a parameter that makes $u = 0$ for the middle of an interval. The set of coefficients a_i , $i = 0, 1, 2, 3, 4$ is different for different segments (Table 1 of Meeus & Simons 2000).

It is important to note that each polynomial is valid only for the period mentioned in the first column of Table 1 and if used for a different period the error could be very large. It should be noted that contrary to the claim of Meeus the largest absolute error is greater than 3 s (for 1627 the error is 3.207471). The error is very large in the beginning and then decreases as one proceeds to higher values of the year. The standard deviation of the error is 0.92546. This large error motivated us to search for a new set of polynomials for which the error does not exceed the 1 s limit. The next section shows that we have been successful in finding such an approximation.

3. New polynomial for Delta T

We have used the same technique as used by Meeus & Simons (2000) introducing a parameter k , which minimizes ‘ u ’ in a specified interval. The only difference is that instead of (2) we have used:

$$u = k + \frac{(\text{year} - 2007)}{100}. \quad (3)$$

Considering polynomials of degree four we have tried to minimize the error to within ± 1 s, selecting a number of segments, i.e., time intervals with an associated polynomial to cover the whole range from AD 1620 to AD 2007. By varying the number of segments and the duration of the segments the best results were obtained for the six segments shown in Table 1 of this work as compared to the eight segments used by Meeus & Simons (2000).

The error analysis has shown that the accuracy is well within ± 1 s for the whole range of the period under consideration. The standard deviation of the error is only 0.3981 which is much better than that in the case of polynomials due to Meeus & Simons (2000). Using this set of new polynomial, the maximum error was found to be 0.990917 s for the year 1806. The major advantage of the polynomials devised by Meeus & Simons is that they give better results for recent years whereas the polynomials devised in this work provide more accurate results in general (absolute error never exceeds 1 s). In the case of Meeus & Simons (2000), the absolute error is more than 1 s in 139 cases out of 381.

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Investigation of the Effects of Expectation Values for Radii on the Determination of Transition Probabilities using WBEPM Theory

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Abstract. Transition probabilities for some excited s–p and p–s transition arrays of neutral nitrogen have been calculated using the weakest bound electron potential model theory (WBEPMT) for the investigation of effects of expectation values of radii. We have used both numerical non-relativistic Hartree-Fock (NRHF) wave functions and numerical Coulomb approximation (NCA) wave functions to calculate expectation values of radii. The transition probability results obtained using the parameters determined with two different wave functions agree well with each other and accepted values taken from NIST for low values of transition probability. However, the NRHF wave functions present better results for p–s transitions, while NCA wave functions are better in s–p transitions for large values of transition probability.

Key words. Neutral nitrogen—atomic data—transition probabilities—WBEPM theory.

1. Introduction

The optical properties of carbon, nitrogen and oxygen atoms are important in both atmospheric and astrophysical applications. Accurate atomic data such as transition probabilities, oscillator strengths and life-times are essential in various fields of astrophysics. Absorption lines observed in sun light related to the transitions between fine structure levels may contain very important information about far stars in galaxy. Moreover, the lighter elements have been proving ground of truly many-electron methodologies. Numerous experimental and theoretical studies have been published over the last 40 years for these elements. Atomic nitrogen is abundant in higher parts of the atmosphere, and its spectrum is apparent in the dayglow. Due to the abundance of atomic nitrogen in stellar atmospheres, spectral lines of NI are prominent features in their absorption spectra, and therefore are becoming increasingly important for all kinds of quantitative analyses (Baclawski *et al.* 2002). The determination of the nitrogen abundance in hot stars is also strongly dependent upon the atomic transition probabilities (Hibbert *et al.* 1991). The accuracy and reliability of radiative transition probabilities for some spectral lines are thus

very important, not only for testing of theoretical models and experimental techniques in atomic systems, but also for possible applications in plasma diagnostics and astrophysical applications. In spite of their extreme importance in atmospheric physics and astrophysics, the multiplet transitions in NI has been examined up until now by applying many theoretical and experimental methods (Beck & Nicolaides 1976).

In this work, in order to investigate the effects of parameters on the transition probabilities, the WBEPM theory has been applied to some excited s–p and p–s transitions in the neutral nitrogen. We employed both NRHF wave functions and NCA wave functions for determination of relevant parameters.

2. Theoretical approximation

The determination of some parameters corresponding to individual energy levels, especially for high lying levels in multi-electron atoms or ions are always a difficult problem for both experimental and theoretical studies. While the calculation procedure for the systems with a few electrons can be carried out easily, the calculations became more difficult and complex in the case of increasing number of electrons. Especially, for the excited states of many-electron atomic or ionic systems, more configurations or basis sets must be considered. Therefore, the calculations became more complicated. Many of the physical properties in multi-electron systems have been studied using some powerful methods such as Multiconfigurational Hartree-Fock (MCHF) (Tong *et al.* 1994; Tachiev & Fischer 2002), Configuration Interaction (CI) (Hibbert *et al.* 1991; Bell *et al.* 1995), Multiconfigurational Dirac-Fock (MCDHF) (Indelicato *et al.* 1988) and R-Matrix method (Seaton 1987; Bell & Berrington 1991). In the well known theoretical methods, atomic wave functions are obtained using many basis-set orbital functions and configurations. The reliability and accuracy of the results depend on the number of basis-set orbital functions chosen. It is difficult to deal with many configuration and orbital basis-set functions and not practicable in calculations. For highly excited states of many electron atoms, it is difficult to obtain appropriate basis-set orbital functions and configurations. Because of the difficulties mentioned above, generally, many of the used methods have considered transitions belong to multiplet levels instead of fine structure transitions and the transitions belong to low lying states rather than highly excited states.

The weakest bound electron potential model theory (WBEPMT) is an efficient method especially for transitions between low excited and highly excited levels. In this method, the determination of some parameters is sufficient for the calculation of the spectroscopic data. The calculations for transition probabilities belong to both low and highly excited levels can be obtained easily and in a shorter time (Zheng *et al.* 2000; Fan & Zheng 2004; Zheng & Wang 2004; Zheng *et al.* 2004). Recently, Çelik *et al.* (2006a) have studied effects of expectation values of radii in some p–d and d–p transitions of neutral nitrogen and this study has shown that changes in the expectation values of radii considerably affect the transition probability results.

The WBEPM theory was proposed by Zheng and he has suggested a new model potential to describe the electronic motion in a multi-electron atom or ion (Zheng 1988).

Zheng *et al.* have employed the WBEPM theory to calculate the energy levels, ionization potentials, transition probabilities, oscillator strengths and lifetimes in many electron atomic or ionic systems. They have obtained very satisfactory physical parameters (Zheng *et al.* 2000a; Fan & Zheng 2004; Zheng & Wang 2004; Zheng *et al.* 2004). The WBEPM theory, based on the idea that electrons in a system can be divided into two groups of electrons; to be the weakest bound electron and non-weakest bound electrons. The weakest bound electron in a given many-electron system is most weakly bound electron to the system compared to the other electrons. By the separation of the electrons in a given system, complex many-electron problem can be simplified as the single electron problem and so can be solved easily (Zheng *et al.* 2004). According to the WBEPM theory, electronic radial wave functions are presented as a function of Laguerre polynomial in terms of some parameters which are determined using the experimental energy data and the expectation values of radii (Zheng *et al.* 2000a),

$$R_{nl}(r) = \left(\frac{2Z^*}{n^*} \right)^{l^*+3/2} \left[\frac{2n^*}{(n-l-1)!} \Gamma(n^* + l^* + 1) \right]^{-1/2} \times \exp \left(-\frac{Z^*r}{n^*} \right) r^{l^*} L_{n-l-1}^{2l^*+1} \left(\frac{2Z^*r}{n^*} \right). \quad (1)$$

Here, Z^* , n^* , l^* quantities are defined to be effective nuclear charge, effective principal quantum number, effective azimuthal quantum number, respectively, and d is an adjustable parameter. Moreover, n^* and l^* parameters have been given to be (Zheng *et al.* 2000a; Zheng & Wang 2002, 2004; Zheng *et al.* 2004; Fan & Zheng 2004):

$$n^* = n + d, \quad l^* = l + d. \quad (2)$$

In the WBEPM theory, these parameters are obtained by solving two equations in equation (3) together.

$$I = -\varepsilon = \frac{Z^{*2}}{2n^{*2}}, \quad \langle r \rangle = \frac{3n^{*2} - l^*(l^* + 1)}{2Z^*}. \quad (3)$$

Here, I is the ionization energy and $\langle r \rangle$ is the expectation value for radius of weakest bound electron. The ionization energies and expectation values for radii of all states must be known for the parameters Z^* , n^* , l^* to be determined. It is well known that some difficulties in obtaining the parameters directly from theory are still present. Therefore, Zheng *et al.* suggest that the values of ionization energy can be taken from the experimental atomic data in the literature and expectation value of radius of the weakest bound electron $\langle r \rangle$ can be obtained by different theoretical methods.

The most important quantity in the calculation of transition probabilities is determination of radial transition integral or radial matrix elements. In this work, we have employed the weakest bound electron potential model theory for determination

Table 1. The parameters obtained using two different wave functions.

Level	n	l	NCA		NRHF		Energy (cm ⁻¹) Ralchenko <i>et al.</i> (2006)
			d	Z^*	d	Z^*	
2p ² (³ P)3s ⁴ P	3	0	-1.165497	1.019477	-1.026685	1.096618	33890.102
2p ² (³ P)3s ⁴ P _{1/2}	3	0	-1.166581	1.019649	-1.025417	1.098157	33941.63
2p ² (³ P)3s ⁴ P _{3/2}	3	0	-1.165871	1.019536	-1.026248	1.097149	33907.87
2p ² (³ P)3s ⁴ P _{5/2}	3	0	-1.164886	1.019380	-1.027399	1.095752	33861.08
2p ² (³ P)3p ⁴ S	3	1	-0.967354	0.878000	-0.987073	0.869483	20474.86
2p ² (³ P)3p ⁴ S _{3/2}	3	1	-0.967354	0.878000	-0.987073	0.869483	20474.86
2p ² (³ P)3p ⁴ P	3	1	-1.057978	0.863903	-0.914685	0.927647	21715.843
2p ² (³ P)3p ⁴ P _{1/2}	3	1	-1.060374	0.863523	-0.912676	0.929278	21750.39
2p ² (³ P)3p ⁴ P _{3/2}	3	1	-1.059100	0.863725	-0.913745	0.928410	21732.01
2p ² (³ P)3p ⁴ P _{5/2}	3	1	-1.056428	0.864149	-0.915981	0.926594	21693.55
2p ² (³ P)3p ⁴ D	3	1	-1.103459	0.856627	-0.942736	0.929222	22387.919
2p ² (³ P)3p ⁴ D _{1/2}	3	1	-1.107858	0.855916	-0.939026	0.932288	22454.82
2p ² (³ P)3p ⁴ D _{3/2}	3	1	-1.106374	0.856156	-0.940279	0.931251	22432.21
2p ² (³ P)3p ⁴ D _{5/2}	3	1	-1.103913	0.856554	-0.942354	0.929538	22394.81
2p ² (³ P)3p ⁴ D _{7/2}	3	1	-1.100550	0.857096	-0.945180	0.927205	22343.88
2p ² (³ P)4s ⁴ P	4	0	-1.134515	1.006234	-1.061183	1.031985	13531.82
2p ² (³ P)4s ⁴ P _{1/2}	4	0	-1.140992	1.006603	-1.054343	1.037111	13603.19
2p ² (³ P)4s ⁴ P _{3/2}	4	0	-1.136948	1.006372	-1.058622	1.033903	13558.54
2p ² (³ P)4s ⁴ P _{5/2}	4	0	-1.130711	1.006020	-1.065172	1.028999	13490.22
2p ² (³ P)4p ⁴ P	4	1	-0.915302	0.940991	-0.837940	0.964590	10211.742

$2p^2(^3P)4p^4P_{1/2}$	4	1	-0.921791	0.940549	-0.831999	0.967985	10245.22
$2p^2(^3P)4p^4P_{3/2}$	4	1	-0.918781	0.940754	-0.834758	0.966408	10229.668
$2p^2(^3P)4p^4P_{5/2}$	4	1	-0.910800	0.941297	-0.842043	0.962247	10188.631
$2p^2(^3P)4p^4D$	4	1	-0.951765	0.938506	-0.866069	0.964891	10402.368
$2p^2(^3P)4p^4D_{1/2}$	4	1	-0.963857	0.937681	-0.854953	0.971315	10466.969
$2p^2(^3P)4p^4D_{3/2}$	4	1	-0.960201	0.937931	-0.858325	0.969365	10447.363
$2p^2(^3P)4p^4D_{5/2}$	4	1	-0.953433	0.938930	-0.864541	0.965773	10411.241
$2p^2(^3P)4p^4D_{7/2}$	4	1	-0.943206	0.939090	-0.873873	0.960390	10357.065
$2p^2(^3P)4p^4S$	4	1	-0.828160	0.946902	-0.909403	0.922648	9780.078
$2p^2(^3P)4p^4S_{3/2}$	4	1	-0.828160	0.946902	-0.909403	0.922648	9780.078
$2p^2(^3P)5s^4P$	5	0	-1.122025	1.003044	-1.078360	1.014338	7341.491
$2p^2(^3P)5s^4P_{1/2}$	5	0	-1.38945	1.003551	-1.060930	1.023828	7413.467
$2p^2(^3P)5s^4P_{3/2}$	5	0	-1.128571	1.003237	-1.071652	1.017987	7369.18
$2p^2(^3P)5s^4P_{5/2}$	5	0	-1.111900	1.002751	-1.088652	1.008747	7299.039
$2p^2(^3P)5p^4P$	5	1	-0.834685	0.967605	-0.817246	0.971656	5921.82
$2p^2(^3P)5p^4P_{1/2}$	5	1	-0.848169	0.967098	-0.804571	0.977254	5954.104
$2p^2(^3P)5p^4P_{3/2}$	5	1	-0.842317	0.967318	-0.810085	0.974817	5940.056
$2p^2(^3P)5p^4P_{5/2}$	5	1	-0.825034	0.967969	-0.826253	0.967686	5898.902
$2p^2(^3P)5p^4D$	5	1	-0.872337	0.966193	-0.840934	0.973544	6012.758
$2p^2(^3P)5p^4D_{1/2}$	5	1	-0.900389	0.965147	-0.814340	0.985405	6082.133
$2p^2(^3P)5p^4D_{3/2}$	5	1	-0.891720	0.965469	-0.822610	0.981711	6060.542
$2p^2(^3P)5p^4D_{5/2}$	5	1	-0.875979	0.966057	-0.837509	0.975068	6021.684
$2p^2(^3P)5p^4D_{7/2}$	5	1	-0.852619	0.966931	-0.859346	0.965363	5964.827

Table 2. Atomic transition probabilities between some excited levels for nitrogen.

Lower state (L)	Upper state (U)	L	U	Wavelength (Å)	Statistical weight		This work (NCA)	This work (NRHF)	NIST values Wiese (2006)
$2s^2 2p^2(^3P)3s$	$2s^2 2p^2(^3P)3p$	$4P$	$4P$	8214.05	12	12	3.21e-01	3.79e-01	3.10e-01[B+]
				8187.111	4	6	9.7e-02	1.14e-01	8.58e-02[B+]
				8190.263	2	4	1.34e-01	1.59e-01	1.27e-01[B+]
				8202.611	2	2	5.3e-02	6.3e-02	4.95e-02[B+]
				8212.98	4	4	4.29e-02	5.05e-02	4.84e-02[B+]
				8218.60	6	6	2.25e-01	2.65e-01	2.23e-01[B+]
				8225.40	4	2	2.67e-01	3.14e-01	2.64e-01[B+]
				8244.66	6	4	1.43e-01	1.68e-01	1.36e-01[B+]
$2s^2 2p^2(^3P)3s$	$2s^2 2p^2(^3P)3p$	$4P$	$4D$	8694.00	12	20	2.74e-01	3.22e-01	2.47e-01[B+]
				8682.666	6	8	2.75e-01	3.23e-01	2.46e-01[B+]
				8721.232	6	6	8.17e-02	9.58e-02	6.75e-02[B+]
				8749.771	6	4	1.34e-02	1.58e-02	1.04e-02[B+]
				8685.788	4	6	1.92e-01	2.26e-01	1.86e-01[B+]
				8714.096	4	4	1.45e-01	1.70e-01	1.28e-01[B+]
				8731.298	4	2	4.51e-02	5.30e-02	3.76e-02[B+]
				8688.535	2	4	1.14e-01	1.34e-01	1.09e-01[B+]
				8705.637	2	2	2.27e-01	2.67e-01	2.10e-01[B+]
$2s^2 2p^2(^3P)3p$	$2s^2 2p^2(^3P)4s$	$4P$	$4P$	12218.9	12	12	6.55e-02	7.52e-02	9.62e-02[B]
				12190.17	6	6	4.5e-02	5.32e-02	7.11e-02[B]
				12292.55	6	4	2.94e-02	3.31e-02	3.76e-02[B]
				12133.29	4	6	1.9e-02	2.31e-02	2.95e-02[B]
				12234.711	4	4	8.7e-03	9.9e-03	1.52e-02[B]
				12301.91	4	2	5.4e-02	6.10e-02	7.41e-02[B]
				12207.26	2	4	2.73e-02	3.13e-02	4.08e-02[B]
				12274.16	2	2	1.09e-02	1.22e-02	1.43e-02[B]

$2s^2 2p^2(^3P)3p$	^4D	$2s^2 2p^2(^3P)4s$	^4P	11291.7 11294.76 11230.16 11316.99 11183.19 11269.29 11326.281 11240.66 11297.35	20	12	1.12e-01 8.95e-02 2.01e-02 7.10e-02 2.24e-03 3.61e-02 5.66e-02 5.64e-03 5.67e-02	1.35e-01 1.08e-01 2.48e-02 8.44e-02 2.19e-03 4.34e-02 6.65e-02 6.83e-03 6.70e-02	1.35e-01[B] 1.08e-01[B] 2.09e-02[B] 9.02e-02[B] 2.01e-03[C+] 4.13e-02[B] 7.36e-02[B] 6.00e-03[B] 6.92e-02[B]
$2s^2 2p^2(^3P)3p$	^4S	$2s^2 2p^2(^3P)4s$	^4P	14402.9 14317.13 14458.55 14552.51	4	12	1.93e-02 1.93e-02 1.92e-02 1.90e-02	1.55e-02 1.58e-02 1.52e-02 1.40e-02	1.24e-02[C] 1.33e-02[C] 1.18e-02[C] 1.10e-02[C]
$2s^2 2p^2(^3P)3p$	^4S	$2s^2 2p^2(^3P)5s$	^4P	7614.19 7589.660 7630.280 7656.151	4	12	7.51e-03 7.6e-03 7.4e-03 7.31e-03	7.05e-03 7.24e-03 6.90e-03 6.73e-03	6.94e-03[D+] 7.01e-03[D+] 6.90e-03[D+] 6.83e-03[D+]
$2s^2 2p^2(^3P)3p$	^4S	$2s^2 2p^2(^3P)6s$	^4P	6294.33 6277.248 6305.658 6323.259	4	12	5.17e-03 5.2e-03 5.0e-03 4.9e-03	5.08e-03 4.71e-03 4.7e-03 4.6e-03	3.72e-03[D+] 3.75e-03[D+] 3.70e-03[D+] 3.67e-03[D+]

(Continued)

Table 2. (Continued)

Lower state (L)	Upper state (U)	L	U	Wavelength (Å)	Statistical weight		This work (NCA)	This work (NRHF)	NIST values Wiese (2006)
$2s^2 2p^2(^3P)4s$	$2s^2 2p^2(^3P)4p$	$4P$	$4D$	31954	12	20	3.24e-02	3.41e-02	3.06e-02[D+]
				31916.7	6	8	3.27e-02	3.42e-02	3.07e-02[D+]
				32478.3	6	6	9.36e-03	9.74e-03	8.74e-03[D+]
				32863.9	6	4	1.51e-03	1.56e-03	1.41e-03[D+]
				31773.3	4	6	2.3e-02	2.42e-02	2.18e-02[D+]
				32142.2	4	4	1.7e-02	1.78e-02	1.60e-02[D+]
				32346.0	4	2	5.22e-03	5.47e-03	4.92e-03[D+]
				31687.4	2	4	1.37e-02	1.45e-02	1.31e-02[D+]
$2s^2 2p^2(^3P)4s$	$2s^2 2p^2(^3P)4p$	$4P$	$4S$	31885.5	2	2	2.70e-02	2.85e-02	2.57e-02[D+]
				26654	12	4	5.06e-02	5.81e-02	4.14e-02[D+]
				26953.1	6	4	2.47e-02	2.81e-02	2.00e-02[D+]
				26465.8	4	4	1.71e-02	1.98e-02	1.41e-02[D+]
				26156.7	2	4	8.75e-03	1.02e-02	7.29e-03[D+]
$2s^2 2p^2(^3P)5s$	$2s^2 2p^2(^3P)5p$	$4P$	$4D$	75260	12	20	8.07e-03	8.31e-03	7.07e-03[D+]
				74950.6	6	8	8.20e-03	8.43e-03	7.15e-03[D+]
				78286.8	6	6	2.2e-03	2.21e-03	1.88e-03[D+]
				80743.0	6	4	3.38e-04	3.35e-04	2.86e-04[D+]
				74211.7	4	6	5.83e-03	6.06e-03	5.16e-03[D+]
				76415.3	4	4	4.12e-03	4.22e-03	3.60e-03[D+]
				77697.2	4	2	1.23e-03	1.25e-03	1.07e-03[D+]
				73913.9	2	4	3.48e-03	3.64e-03	3.11e-03[D+]
$2s^2 2p^2(^3P)5s$	$2s^2 2p^2(^3P)5p$	$4P$	$4D$	75112.6	2	2	6.68e-03	6.94e-03	5.92e-03[D+]

Note: Estimated accuracy rates from NIST: B+ ≤ 7%, B ≤ 10%, C+ ≤ 18%, D+ ≤ 40%.

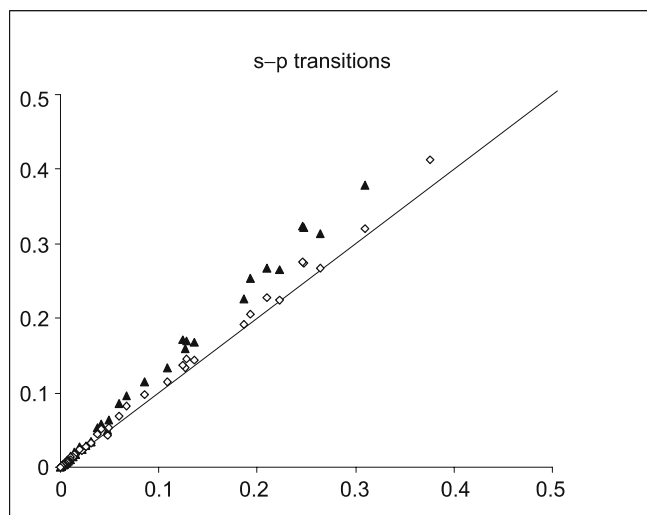
of radial transition integrals. According to the WBEPM theory, the radial transition integral for transitions between two excited levels is given as (Zheng *et al.* 2000a),

$$\begin{aligned}
 \langle n_i, l_i | r^k | n_f, l_f \rangle &= \int_0^\infty r^{k+2} R_{n_i l_i}(r) R_{n_f l_f}(r) dr \\
 &= (-1)^{n_f + n_i + l_f + l_i} \left(\frac{2Z_f^*}{n_f^*} \right)^{l_f^*} \left(\frac{2Z_i^*}{n_i^*} \right)^{l_i^*} \times \left(\frac{Z_f^*}{n_f^*} - \frac{Z_i^*}{n_i^*} \right)^{-l_f^* - l_i^* - k - 3} x \\
 &\quad \times \left[\frac{n_f^{*4} \Gamma(n_f^* + l_f^* + 1)}{4Z_f^{*3} (n_f - l_f - 1)!} \right]^{-1/2} x \left[\frac{n_i^{*4} \Gamma(n_i^* + l_i^* + 1)}{4Z_i^{*3} (n_i - l_i - 1)!} \right]^{-1/2} \\
 &\quad \times \sum_{m_1=0}^{n_f - l_f - 1} \sum_{m_2=0}^{n_i - l_i - 1} \frac{(-1)^{m_2}}{m_1! m_2!} \left(\frac{Z_f^*}{n_f^*} - \frac{Z_i^*}{n_i^*} \right)^{m_1 + m_2} \times \left(\frac{Z_f^*}{n_f^*} + \frac{Z_i^*}{n_i^*} \right)^{-m_1 - m_2} x \\
 &\quad \times \Gamma(l_f^* + l_i^* + m_1 + m_2 + k + 3) \times \sum_{m_3=0}^S \left(\frac{l_i^* - l_f^* + k + m_2 + 1}{n_f^* - l_f^* - 1 - m_1 - m_3} \right) \\
 &\quad \times \left(\frac{l_f^* - l_i^* + k + m_1 + 1}{n_i^* - l_i^* - 1 - m_2 - m_3} \right) \times \binom{l_i^* + l_f^* + k + m_1 + m_2 + m_3 + 2}{m_3}, \quad (4)
 \end{aligned}$$

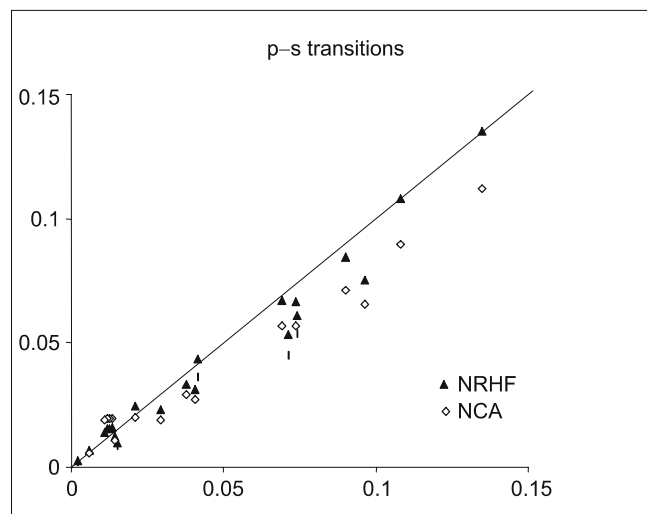
where $S = \min\{n_f - l_f - 1 - m_1, n_i - l_i - 1 - m_2\}$ and $k > -l_f^* - l_i^* - 3$. In this work, expectation values of radii for all excited states have been calculated both by NCA wave functions (Lindgard & Nielsen 1977) and by numerical NRHF wave functions (Gaigalas & Fischer 1996) in determination of parameters required for the transition probability calculations. The necessary experimental energy values have been taken from NIST (Ralchenko *et al.* 2006). The parameters obtained by using the two different wave functions are presented in Table 1. Then, making use of these parameters, transition probabilities have been calculated in neutral nitrogen.

3. Results and conclusions

In order to investigate the effects of expectation values for radii, the transition probabilities have been calculated for some s-p and p-s transitions of excited atomic nitrogen in the WBEPM theory framework. The results obtained in this work are presented in Table 2. Expectation values of radii for all states have been calculated by using both the NCA wave functions and the numerical NRHF wave functions to determine the Z^* and d parameters required for the calculation of the transition probabilities. The results of transition probability have been compared with each other and accepted values are taken from NIST (Wiese 2006) which contains recommended values for many transitions. The NIST values are given together with their accuracy rating in relevant columns of tables. The estimated uncertainties of accepted values are $\pm 7\%$ for 3s-3p transitions, ± 10 -40% for 3p-4s, 5s, 6s transitions and $\pm 40\%$ for



(a)



(b)

Figure 1. Transition probability results calculated with parameters obtained using NRHF and NCA wave functions *versus* accepted values taken from NIST for neutral nitrogen. (a) s–p transitions (b) p–s transitions.

4s–4p and 5s–5p. It can be seen from Table 2 that there is an agreement between data obtained in the present work and the accepted values from NIST. Experimental and theoretical data for comparison are quite limited for some excited transitions. On the other hand, some data have not been sensitively tested for both the multiplet values and the individual lines, especially in transitions belonging to highly excited levels. Therefore, the results obtained from our calculations have been compared to only

NIST data. It is well known from literature that the WBEPM theory is an efficient method especially for transitions between excited and highly excited levels. Accuracy and reliability of the spectroscopic data results obtained from the WBEPM theory strongly depend on the accuracy and reliability of the expectation values of radii and experimental ionization energies which are used to determine Z^* , n^* and l^* parameters. Since experimental ionization energy data are very precise, the determination of the expectation values of the radii is more crucial than energy values in the WBEPM theory. Zheng *et al.* have employed only NCA method in all the calculations carried out using the WBEPM theory for the expectation values of radii. In this study, we have employed more sophisticated numerical NRHF wave functions than the NCA wave functions which are commonly employed in the traditional WBEPM theory procedure. Previously, we employed numerical NRHF wave functions for the determination of relevant parameters in atomic nitrogen (Çelik *et al.* 2006a, 2007), lithium (Çelik 2007) and fluorine (Çelik *et al.* 2006b) and obtained very satisfactory transition probability and oscillator strength results by using the WBEPM theory.

In order to observe the effects of expectation values for radii, we have plotted transition probability results calculated with parameters obtained using NRHF and NCA wave functions *versus* accepted values taken from NIST in both some s–p transitions and p–s transitions. It can be seen clearly from Table 2 and Fig. 1 that our results are in agreement with the accepted values taken from NIST. It can be concluded that, according to Fig. 1, our results obtained by using both NRHF wave functions and NCA wave functions for low values of transition probability are in agreement with NIST data. The parameters obtained from NRHF wave functions present better results in p–s transitions, while the parameters obtained from NCA wave functions present better results in s–p transitions for large values of transition probability.

The obtained results indicate that accuracy of transition probabilities in the WBEPM theory framework strongly depends on the accuracy and reliability of the expectation values for radii. In the WBEPM theory, although relativistic effects are neglected except for the observed binding energies, many of the spectroscopic data results are in very good agreement with the results obtained from theoretical methods and experimental measurements in the literature. By courtesy of this method, physical parameters can be calculated for both highly excited states and low lying states without any increase of complexity in calculation process. The transitions between individual lines in both low and highly excited levels have a great importance in many areas of the physics, especially in astrophysics. The calculations of spectroscopic data will be better, if the expectation values of radii belonging to the levels are accurate enough.

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Gravitation Field Dynamics in Jeans Theory

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Abstract. Closed system of time equations for nonrelativistic gravitation field and hydrodynamic medium was obtained by taking into account binary correlations of the field, which is the generalization of Jeans theory. Distribution function of the system was built on the basis of the Bogolyubov reduced description method. Calculations were carried out up to the first order of a perturbation theory in interaction. Adiabatic and entropic types of perturbations were corrected and two new types of perturbations were found.

Key words. Time equations—Jeans theory—Bogolyubov reduced description method—perturbation types.

Jeans theory is the fundamental basis for self-gravitating gas study (Zeldovich & Novikov 1983) and is known for the simplicity of the model. Earlier generalizations of the Jeans theory for more complicated gas models are actively studied, for example, the model with Burnett terms (Garcia-Colin & Sandoval-Villalbazo 2005), magnetization (Lee & Hong 2007), heat flux (Sandoval-Villalbazo & Garcia-Perciante 2007), binary systems (Tsiklauri 1998). In this paper, on the other hand, generalization was built by taking into account freedom degrees of nonrelativistic longitudinal gravitation field. New independent variables (second correlations of the field) are introduced. This gives us the possibility to estimate the influence of the mentioned field on perturbation motion in the gas. This problem is analogous to the Coulomb electronic plasma with Langmuir (plasma) waves (Sokolovsky & Stupka 2006).

1. Hamilton function

We will begin with nonrelativistic action for the system of massive particles and Newtonian field (Landau & Lifshitz 1980)

$$S = \iiint \left(\frac{\rho v^2}{2} - \rho \phi - \frac{(\nabla \phi)^2}{8\pi G} \right) dV dt. \quad (1)$$

Let us make the replacement of generalized field co-ordinate

$$\varphi = \partial_t \lambda. \quad (2)$$

If we consider equation (1) as nonrelativistic limit of general relativity action, then $\varphi = -c^2 h_{00}/2$, where h_{00} is a small addition term to the proper component of Galilei metrics.

The co-ordinates shift by small four-vector ξ_α (the Greek indices are 0, 1, 2, 3) causes change of metric tensor (Landau & Lifshitz 1980)

$$h'_{\alpha\beta} = h_{\alpha\beta} - \frac{\partial \xi_\alpha}{\partial x^\beta} - \frac{\partial \xi_\beta}{\partial x^\alpha},$$

by comparing with equation (2) one can see $\xi_0 = \lambda/c$. This means that equation (2) corresponds to returning from the co-ordinate system with $h_{00} = 0$ to the laboratory system.

Let us denote $\mathbf{A} = \nabla\lambda$, where the components of $\mathbf{A}$ coincide with $-ch_{0i}$ of the system with $h_{00} = 0$ at $\xi_i = 0$: $-ch_{0i} = \nabla_i\lambda$. In a linear theory the gravitation vector potential is formed from the components h_{0i} . Similarly, one could name the vector $\mathbf{A}$ as longitudinal vector potential, however, one should remember that it is a new generalized co-ordinate of the field exactly in the laboratory system.

Then, the second term in equation (1) integrating by parts over time and using equation of continuity:

$$\partial_t \rho + \nabla j = 0, \quad (3)$$

and after integrating by parts over space, and in the third term simply changing the order of derivatives we obtain:

$$S = \iiint \left(\frac{\rho v^2}{2} - j\mathbf{A} - \frac{(\partial_t \mathbf{A})^2}{8\pi G} \right) dV dt. \quad (4)$$

In equation (4), unlike equation (1), the generalized co-ordinate of gravitation field is a dynamic one and it is easy to build the Hamiltonian function by standard procedure (Sokolovsky & Stupka 2006). Let us introduce the correspondent velocity $\mathbf{E} = \partial_t \mathbf{A}$, and nonrelativistic Hamilton function of the system has the form:

$$\begin{aligned} \hat{H} &= \hat{H}_0 + \hat{H}_{\text{int}} = (\hat{H}_m + \hat{H}_f) + (\hat{V}_1 + \hat{V}_2), \\ \hat{H}_f &= - \int d^3x \frac{E(x)^2}{8\pi}, \\ \hat{V}_1 &= - \int d^3x \hat{\mathbf{A}}_n(\mathbf{x}) \hat{j}_{\text{on}}(\mathbf{x}), \\ \hat{V}_2 &= \frac{1}{2} \int d^3x \hat{\mathbf{A}}^2(\mathbf{x}) \hat{\rho}(\mathbf{x}); \\ \hat{j}_{\text{on}}(\mathbf{x}) &= \sum_a \hat{j}_{\text{on}a}(\mathbf{x}) = \hat{\pi}_{\text{on}}(\mathbf{x}), \end{aligned} \quad (5)$$

where $\hat{H}_m$ and $\hat{H}_f$ are Hamilton functions for free medium particles and longitudinal gravitation field respectively ($\text{rot } E = 0$). The last expression (5) contains free (i.e., without interaction) momentum density $\hat{\pi}_{\text{on}}(\mathbf{x})$ (such values have additional subscript o), that involve a question about gauge invariance, which will be solved in a standard way (Sokolovsky & Stupka 2006; Landau & Lifshitz 1980).

2. Distribution function

The considered system will be described by field strength $\eta_{1nx} \equiv \mathbf{E}_n(\mathbf{x}, t)$, its longitudinal vector potential $\eta_{2nx} \equiv \mathbf{A}_n(\mathbf{x}, t)$, binary correlations $(\hat{\mathbf{E}}_n(x) \hat{\mathbf{E}}_l(x'))^t$ (we denote them as variables $\eta_\alpha(t)$) and by densities of energy $\zeta_0(x, t) \equiv \varepsilon(\mathbf{x}, t)$, momentum $\zeta_n(\mathbf{x}, t) \equiv \boldsymbol{\pi}_n(\mathbf{x}, t)$, mass $\zeta_4(x, t) \equiv \rho(\mathbf{x}, t)$ of gas subsystem (variables $\zeta_\mu(\mathbf{x}, t)$). A similar parameter set was used in Sokolovsky & Stupka (2006). Gauge invariant densities of momentum and energy can be introduced in the usual manner. For example, gauge invariant mass velocity u_{an} of the electron subsystem is given by the expression:

$$\mathbf{u}_n(\mathbf{x}, t) = \mathbf{u}_{\text{on}}(\mathbf{x}, t) - \mathbf{A}_n(\mathbf{x}, t) \rho(\mathbf{x}, t), \quad (6)$$

where u_{on} is free velocity. This allows us to use the Galilei transformation and to obtain results in a gauge invariant form.

Equations of motion for microscopic values of reduced description parameters in terms of the gauge invariant densities have the standard form:

$$\begin{aligned} \hat{\mathbf{E}}_n(\mathbf{x}) &= 4\pi G \hat{\mathbf{j}}_n(\mathbf{x}), \quad \hat{\mathbf{A}}_n(\mathbf{x}) = -\hat{\mathbf{E}}_n(\mathbf{x}); \\ \hat{\rho}(\mathbf{x}) &= -\frac{\partial \hat{j}_l(x)}{\partial x_l}, \quad \hat{\varepsilon}(\mathbf{x}) = -\frac{\partial \hat{q}_n(x)}{\partial x_n} + \hat{j}_n(x) \hat{\mathbf{E}}_n(\mathbf{x}), \\ \hat{\pi}_l(x) &= -\frac{\partial \hat{t}_{ln}(x)}{\partial x_n} + \hat{\rho}(x) \hat{\mathbf{E}}_l(x). \end{aligned} \quad (7)$$

These formulae allow us to write the corresponding microscopic equations for products of values $\hat{\mathbf{E}}_n(\mathbf{x})$. Averaging such equations with the nonequilibrium distribution function of the system $\rho(\eta(t), \zeta_o(t))$ we obtain a closed system of equations for parameters describing the system's state (it is convenient to do this by using non-gauge invariant medium variables ζ_o).

To construct distribution function $\rho(\eta(t), \zeta_o(t))$ for this case, one can use the Bogolyubov reduced description method (Akhiezer & Peletminsky 1981) starting from the Liouville equation

$$\partial_t \rho(\eta(t), \zeta_o(t)) = \{\hat{H}, \rho(\eta(t), \zeta_o(t))\} \equiv \mathbf{L} \rho(\eta(t), \zeta_o(t)). \quad (8)$$

Field microscopic values and their correlations satisfy the Peletminsky–Yatsenko condition $\mathbf{L}_m \hat{\eta}_\alpha = -i \sum_{\alpha'} c_{\alpha\alpha'} \hat{\eta}_{\alpha'}$. Using a boundary condition of complete correlation weakening, according to Akhiezer & Peletminsky (1981) and Sokolovsky & Stupka (2005), we obtain the following integral equation for distribution function $\rho(\eta, \zeta_o)$:

$$\begin{aligned} \rho(\eta, \zeta_o) &= \rho_q(\eta,) w(\zeta_o) + \int_0^{+\infty} d\tau e^{\tau \mathbf{L}_m} \left\{ e^{\tau \mathbf{L}_f} \left(\mathbf{L}_{\text{int}} \rho(\eta, \zeta_o) + 4\pi G \sum_\alpha \frac{\partial \rho(\eta, \zeta_o)}{\partial \eta_\alpha} \right. \right. \\ &\quad \left. \left. \times j_\alpha(\eta, \zeta_o) - \sum_\mu \int d\mathbf{x} \frac{\partial \rho(\eta, \zeta_o)}{\partial \zeta_{\mu o}(\mathbf{x})} M_\mu(\mathbf{x}, \eta, \zeta_o) \right) \right. \\ &\quad \left. - \rho_q(\eta) \mathbf{L}_m w(\zeta_o) \right\} \Big|_{\eta \rightarrow e^{i\tau c\eta}} \end{aligned} \quad (9)$$

Further, it is convenient to consider the equation (9) in the picture of spatial dependence of microscopic values by Baryakhtar–Peletminsky (Akhiezer & Peletminsky

1981) using distribution function $\rho(\mathbf{x}, \eta, \zeta_o)$ instead of $\rho(\eta, \zeta_o)$. Then one can choose $w(\mathbf{x}, \zeta_o)$ in the form of local in $\zeta_{o\mu}(\mathbf{x})$ quasi-equilibrium distribution function of the medium and $\rho_q(\eta)$ as arbitrary distribution function of the field. Functions $j_\alpha(\eta, \zeta_o)$, $M_\mu(\mathbf{x}, \eta, \zeta_o)$ are defined by equations of motion for parameters describing the system

$$\begin{aligned}\dot{\eta}_\alpha(t) &= i \sum_{\alpha'} c_{\alpha\alpha'} \eta_{\alpha'}(t) + 4\pi G j_\alpha(\eta(t), \zeta_o(t)), \\ \dot{\zeta}_{o\mu}(x, t) &= M_\mu(x, \eta(t), \zeta_o(t)).\end{aligned}\quad (10)$$

In the Baryakhtar–Peletminsky picture equation (9) is solvable in a perturbation theory in small interaction ($\hat{V}_1 \sim \lambda^1$, $\hat{V}_2 \sim \lambda^2$, $\lambda \ll 1$). The weakness of EM interaction allows us to build the perturbation theory in small parameter $\lambda = \Omega/\omega_0$, where Ω , ω_0 are Jeans and collision frequencies starting from estimations $L_s \sim \lambda^0$, $L_{bo} \sim \lambda^0$, $L_1 \sim \lambda^1$, $L_2 \sim \lambda^2$ for the contribution in Liouville operator (Sokolovsky & Stupka 2005).

It was found that in the local rest reference system of the gas subsystem:

$$\begin{aligned}\rho^{0(0)}(x, \eta, \zeta_o) &= \rho_q(\eta) w(\zeta_o(x)), \\ w(\zeta_o) &= \exp \beta \{ \Omega(\beta, \mu) - \hat{H}_m + \mu \hat{M} \}; \\ \rho^{0(1)}(x, \eta, \zeta_o) &= \int_{-\infty}^0 d\tau \int dx' \sum_{\alpha} v_{n\alpha}(x', \tau) \\ &\quad \times \left(\{ \rho_q w(\zeta_o(\mathbf{x})), \hat{\eta}_\alpha(\hat{j}_{\text{on}}(\mathbf{x}' - \mathbf{x}, \tau) + \hat{\rho}(\mathbf{x}' - \mathbf{x}, \tau) \mathbf{u}_n(x)) \} \right. \\ &\quad + \rho_q \hat{\eta}_\alpha \sum_{\mu} \frac{\partial w(\zeta_o(\mathbf{x}))}{\partial \zeta_{o\mu}} \text{Sp}_m w(\zeta_o(\mathbf{x})) \{ \hat{\zeta}_\mu(0), (\hat{j}_{\text{on}}(\mathbf{x}' - \mathbf{x}) \\ &\quad \left. + \hat{\rho}(\mathbf{x}' - \mathbf{x}) \mathbf{u}_n(x)) \} + w(\zeta_o(\mathbf{x})) \hat{\rho}(\mathbf{x}') u_n(x') \sum_{\alpha'} \frac{\partial \rho_q}{\partial \eta_{\alpha'}} \varepsilon_{\alpha'\alpha} \right),\end{aligned}\quad (11)$$

where $\hat{A}_n(x)$ and $\hat{j}_{\text{an}}(\mathbf{x})$ in the Dirac picture and matrix $\varepsilon_{\alpha\alpha'}$ were introduced

$$\begin{aligned}\hat{A}_n(x, \tau) &\equiv e^{-\tau \mathbf{L}_f} \hat{A}_n(x) = \sum_{\alpha} \mathbf{v}_{n\alpha}(x, \tau) \hat{\eta}_\alpha, \\ \hat{j}_{\text{on}}(x, \tau) &= e^{-\tau \mathbf{L}_m} \hat{j}_{\text{on}}(\mathbf{x}), \quad \varepsilon_{\alpha\alpha'} \equiv i \text{Sp}_f \rho_q \{ \hat{\eta}_\alpha, \hat{\eta}_{\alpha'} \}\end{aligned}\quad (12)$$

($\mathbf{v}_{n\alpha}(\mathbf{x}, \tau)$ are known functions), w is equilibrium distribution function of the gas subsystem, $\hat{H}_m$, $\hat{M}$ are the Hamiltonian and mass density of the gas subsystem. Zero superscript in $\rho^{0(0)}(x, \eta, \zeta)$, $\rho^{0(1)}(x, \eta, \zeta)$ and further indicates, that the corresponding value belongs to the system of the local rest. The first term in equation (11) contains

the expression:

$$\begin{aligned} & \int_{-\infty}^0 d\tau \int d\mathbf{x} \{ \hat{\mathbf{A}}_n(\mathbf{x}, \tau) \hat{\mathbf{j}}_n(\mathbf{x}, \tau), \rho_q(\eta) w \} \\ &= \int_{-\infty}^0 d\tau \int d\mathbf{x} (\{ \hat{\mathbf{A}}_n(\mathbf{x}, \tau), \rho_q(\eta) \} \hat{\mathbf{j}}_n(\mathbf{x}, \tau) w + \rho_q(\eta) \hat{\mathbf{A}}_n(\mathbf{x}, \tau) \{ \hat{\mathbf{j}}_n(\mathbf{x}, \tau), w \}), \end{aligned} \quad (13)$$

in which the last Poisson brackets can be rewritten in the form as in equation (5) (Sokolovsky & Stupka 2005)

$$\{ \hat{\mathbf{j}}_n(\mathbf{x}, \tau), w \} = -\beta \{ \hat{\mathbf{j}}_n(\mathbf{x}, \tau), \hat{H}_m \} w = \beta \frac{\partial \hat{\mathbf{j}}_n(\mathbf{x}, \tau)}{\partial \tau} w \quad (14)$$

($\beta = T^{-1}$). Then integrating by parts and taking into account the boundary condition of the complete correlation weakening, we see that the lower integral limit disappears and equation (11) takes the form

$$\begin{aligned} \rho^0(\mathbf{x}, \eta, \zeta) &= \rho_q(\eta) w - \int_{-\infty}^0 d\tau \int d\mathbf{x} \left(\{ \hat{\mathbf{A}}_n(\mathbf{x}, \tau), \rho_q(\eta) \} \hat{\mathbf{j}}_n(\mathbf{x}, \tau) w \right. \\ &\quad \left. + w(\zeta_o(\mathbf{x})) \sum_{\alpha} \mathbf{v}_{n\alpha}(\mathbf{x}', \tau) \hat{\rho}(\mathbf{x}') \mathbf{u}_n(\mathbf{x}') \sum_{\alpha'} \frac{\partial \rho_q}{\partial \eta_{\alpha'}} \varepsilon_{\alpha'\alpha} \right) \\ &\quad - \beta \int_{-\infty}^0 d\tau \int d\mathbf{x} \rho_q(\eta) \hat{\mathbf{E}}_n(\mathbf{x}, \tau) \hat{\mathbf{j}}_n(\mathbf{x}, \tau) w \\ &\quad - \beta \int d\mathbf{x} \rho_q(\eta) \hat{\mathbf{A}}_n(\mathbf{x}) \hat{\mathbf{j}}_n(\mathbf{x}) w. \end{aligned} \quad (15)$$

The last term in equation (15) will be dismissed with term from $\hat{V}_2$.

3. Equations of motion and perturbations

To simplify further discussion we will keep in equations, terms up to the first order in λ .

$$\begin{aligned} \frac{\partial \rho}{\partial t} &= -\frac{\partial \rho u_n}{\partial x_n}, \\ \frac{\partial \mathbf{u}_l}{\partial t} &= -\mathbf{u}_n \frac{\partial \mathbf{u}_l}{\partial x_n} - \frac{1}{\rho} \frac{\partial (p + p^{(1)})}{\partial x_l} + \mathbf{E}_l + \frac{1}{\rho} (\hat{\rho} \hat{\mathbf{E}}_l)^0, \\ \frac{\partial \varepsilon^0}{\partial t} &= -\frac{\partial (\varepsilon^0 u_n + q_n^{(0(1))})}{\partial x_n} - p \frac{\partial \mathbf{u}_l}{\partial x_l} + \rho \mathbf{u}_n \mathbf{E}_n + (\hat{\mathbf{j}}_n \hat{\mathbf{E}}_n)^0. \end{aligned} \quad (16)$$

Here $p, p^{(1)}$ are the local equilibrium and the first order contribution to pressure of the gas subsystem; $q_n^{(0(1))}$ is the first order contribution to energy flux of the gas subsystem.

The medium-field correlations are given by the expressions:

$$\begin{aligned}
 (\hat{\mathbf{j}}_n \hat{\mathbf{E}}_l) &= (\hat{\mathbf{j}}_n \hat{\mathbf{E}}_l)^0 + (\hat{\rho} \hat{\mathbf{E}}_l) \mathbf{u}_n, \\
 (\hat{\zeta}_{o\mu n} \hat{\mathbf{E}}_l) &= \lambda_{\mu n, \alpha}(\mathbf{x}, \zeta(x)) (\hat{\eta}_\alpha \hat{\mathbf{E}}_l) + S_{\mu n, 1l}(\mathbf{x}, \zeta(x)); \\
 S_{\mu n, \alpha}(\mathbf{x}, \zeta) &= 8\pi \int_{-\infty}^0 d\tau \int d\mathbf{x}' \sum_{\alpha'} \varepsilon_{\alpha\alpha'} \mathbf{v}_{l\alpha'}(\mathbf{x}', \tau) \\
 &\quad \times \text{Sp}_m w(\zeta) (\hat{\mathbf{j}}_{\text{ol}}(\mathbf{x}', \tau) + \hat{\rho}(\mathbf{x}', \tau) \mathbf{u}_l - \rho \mathbf{u}_l) \hat{\zeta}_{o\mu n}(\mathbf{x}), \\
 \lambda_{a\mu n, \alpha}(\mathbf{x}, \zeta) &= -\beta \int_{-\infty}^0 d\tau \int d\mathbf{x}' v_{l\alpha}(\mathbf{x}') \text{Sp}_m w^o(\zeta^o) \hat{\zeta}_{oa\mu n}(\mathbf{x}) \\
 &\quad \times \left(\sum_b \hat{\mathbf{j}}_{\text{obl}}(\mathbf{x}', \tau) + \hat{\rho}(\mathbf{x}', \tau) \mathbf{u}_l \right). \tag{17}
 \end{aligned}$$

The first order contribution to hydrodynamic fluxes has the form $\zeta_{oa\mu n}^{(1)} = \lambda_{a\mu n, 1l}(\mathbf{x}, \zeta(x)) \mathbf{E}_l$, where function $\lambda_{a\mu n, \alpha}$ is defined in equation (17). In the considered approximation, relations (11) and (14) give the following equations for average gravitation field and correlations of the field

$$\frac{\partial \mathbf{E}_n}{\partial t} = 4\pi G \rho u_n, \quad \frac{\partial (\hat{\mathbf{E}}_n \hat{\mathbf{E}}_l)}{\partial t} = 4\pi G (\hat{\mathbf{j}}_n \hat{\mathbf{E}}_l). \tag{18}$$

From equation (17), it follows that we received a closed system of equations (16) and (18) for gravitation field and gas. These equations can be applied for studying field dynamics and correlation phenomena in the system. Suppose the correlations have their equilibrium value $(EE)_{nn} = 4\pi T$, one obtains equations of Jeans theory, but the first equation (18) allows zero solution, which was forbidden by Poisson equation.

Further, we will ignore spatial dispersion and consider zero correlation radius approximation using notation of a kind $(\hat{\mathbf{E}}_n(\mathbf{x}) \hat{\mathbf{E}}_l(\mathbf{x}')) = (\hat{E} \hat{E})_{nl}(\mathbf{x}) \delta(\mathbf{x} - \mathbf{x}')$. Linearization of equations (16)–(18) near the equilibrium after the Fourier transformation gives the following equations for deviations of the corresponding values from the equilibrium:

$$\begin{aligned}
 \frac{\partial \rho}{\partial t} &= -i\rho_0 k u, \quad \frac{\partial u}{\partial t} = -\frac{i}{\rho_0} k(p_T T + p_\rho \rho) + E, \\
 \varepsilon_T^0 \frac{\partial T}{\partial t} &= -ik\lambda_4 E - (w - \varepsilon_\rho^0 \rho_0) iku + \lambda_2 \{(EE)_{nn} - 4\pi T\}; \tag{19}
 \end{aligned}$$

$$\frac{\partial E}{\partial t} = 4\pi G \rho_0 u, \quad \frac{\partial (EE)_{nn}}{\partial t} = 4\pi \lambda_1 \{(EE)_{nn} - 4\pi T\}, \tag{20}$$

(we do not use a special notation for the deviations; ρ_0 is equilibrium mass density of electrons). Here only longitudinal parts of velocity vector fields were kept, constant coefficients $\lambda_1 \delta_{nl} = -G \int d\mathbf{x} \lambda_{4n, 1l}(\mathbf{x}, \zeta)$, $\lambda_2 \delta_{nl} = \lambda_{4n, 1l}(\mathbf{x} = 0, \zeta)$,

$\lambda_4 \delta_{nl} = \int d\mathbf{x} \lambda_{0n,1l}(x, \zeta)$ were introduced and notations $p = p_T T + p_\rho \rho$, $\varepsilon^0 = \varepsilon_T^0 T + \varepsilon_\rho^0 \rho$ were used. So, we have five equations (19) and (20), which form a closed system. It allows us to study the influence of average field and field correlation degrees of freedom on such systems. And we obtain two new perturbation types: field one, that characterized by λ_4 and correlation one, that characterized by λ_1 and λ_2 . To solve the system we assume the time-dependence to have the form $\sim \exp(\omega t)$. After some calculation we obtain the following corrected dispersion law for the known adiabatic types (Zeldovich & Novikov 1983) for small wave vector and small interaction:

$$\omega = \pm (\Omega + O(\lambda^2)) + \left(\mp \frac{u^2}{2\Omega} - \frac{\Delta}{2} + \frac{\psi\theta}{2\Omega^2} + O(\lambda^2) \right) k^2 + O(k^3), \quad (21)$$

where $\Omega = \sqrt{4\pi G m}$ is Jeans frequency, $\Delta \equiv 4\pi \lambda_4 p_T / \varepsilon_T^0$, $\psi \equiv p_T(\varepsilon^0 + p - \varepsilon_\rho^0 \rho) / \rho \varepsilon_T^0$, $\theta \equiv 4\pi \lambda_2 / \varepsilon_T^0$ and $u = \sqrt{p_\rho + \psi}$ is the sound velocity. For the field type $\omega = 0$ and for entropic and correlation types:

$$\omega = \left(\frac{\vartheta - \theta \pm |\vartheta - \theta|}{2} + O(\lambda^2) \right) \pm \left(\frac{\psi\theta(\vartheta - \theta \pm |\vartheta - \theta|) + \Delta(\vartheta + \theta \pm |\vartheta - \theta|)\Omega^2}{2\Omega^2|\vartheta - \theta|} + O(\lambda^2) \right) k^2 + O(k^3), \quad (22)$$

where $\vartheta = 4\pi \lambda_1$. Together with the sound velocity the expression (21) contains new terms connected with average field and field correlations because kinetic coefficients λ_4 and λ_2 , which can be a leading contribution compared with an ordinary expression (Zeldovich & Novikov 1983). Expression (22) gives an essentially new effect due to new kinetic coefficients λ_4 , λ_1 and λ_2 .

4. Conclusions

We have generalized Jeans theory by taking into account binary correlations of non-relativistic gravitation field and using time equation for the average field. This has been done by using Bogolyubov reduced description method and mass conservation law starting from action for nonrelativistic gas and Newtonian field. As a result, the solution of linearized system has been proposed, that corrected known adiabatic perturbation eigenvalues and introduced two new ones, one of which is zero, but the other is coupled with corrected entropic perturbation. This, for example, changes the Jeans mass and may be useful in galaxy organization theories.

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Observational Evidences for Multi-component Magnetic Field Structure in Solar Flares

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Abstract. Two solar flares of 25 July 1981 and 5 November 2004 of importance 2N and M4.1/1B, respectively, were investigated using observational data obtained with the Echelle spectrograph of the Kyiv University Astronomical Observatory. Stokes I and V profiles of the FeI lines 5233, 5247.1, 5250.2, 5250.6, 5576.1 and of CrI 5247.6 Å have been analyzed. We found several evidences for the existence of spatially unresolved magnetic field structures with kG strengths. In particular, the values of the measured average longitudinal field $B_{\parallel}$ depend on the Lande factors g of the lines: in general, $B_{\parallel}$ increases with increasing factor g . Analogously, the observed line ratio $B_{\parallel}(5250.2)/B_{\parallel}(5247.1)$ is increasing with increasing distance $\Delta\lambda$ from the line center. The observed Stokes V profiles show some deviations from that of an assumed homogeneous field, presented by the Stokes I gradient, $dI/d\lambda$. A comparison with the non-split line FeI 5576.1 Å shows that some of these deviations are real and indicate the presence of subtelescopic magnetic elements with discrete field strengths of several kG. The lines with large Lande factors have considerable broadenings of the Stokes I profiles, indicating a strong background magnetic field of mixed polarity. On the basis of all these data we conclude that a four-component magnetic field structure is a possible explanation. The field strengths are about ± 1.05 kG in the background field, and 1.3–1.5, 3.9–4.0, and 7.4–7.8 kG at level of middle photosphere ($h \approx 300$ km) in the spatially unresolved, small-scale magnetic elements.

Key words. Solar flares—small-scale magnetic fields—multi-component structure.

1. Introduction

Solar flares are closely connected to magnetic fields and rapid reconnection of magnetic fields plays a key role in these events (see, e.g., Severny 1988; Priest 1986). However, Parker (2001) emphasized that “we do not understand why the magnetic field is in the intensely fibril state that appears at the surface”, and “the conditions for

reconnection between fibrils necessarily becomes involved with the dynamical state of the fibrils themselves, about which little is known . . .”. Really, detailed magnetic characteristics of the flare volume are practically unknown due to essential problems with the magnetic field measurements in solar active regions and flares. This problem has been discussed by many authors, e.g., by Lozitska & Lozitsky (1982), Moore *et al.* (1984), and Staude & Hofmann (1988). It is due to the strong dependence of some parameters of the used magneto-sensitive spectral lines on both magnetic field structure and thermodynamic conditions. Some types of instruments, e.g., the traditional magnetographs of the Babcock (1953) type, are practically useless to measure reliably the magnetic field in flare, in the case of strong flare emission with essential profile changes in the line core in particular. On the other hand, instruments such as Fourier transform spectrometers (Brault 1978) are more suitable for this case, but they are too slow to investigate such rapid phenomena as flares. A successful investigation of a flare using such an instrument was made, but in one spectral line only (Deming *et al.* 1990). However, data of many spectral lines are needed to determine the true magnetic structure in flares. Traditional spectral observations using analyzers for circularly polarized light are also suitable for this purpose. Though such data are inferior to those of a Fourier transform spectrometer in sensitivity and precision, they allow faster recording of a wide spectral range, which is of basic importance for flares.

There exist spectral observations pointing out the small-scale magnetic field structure in flares. A two-component flare emission in metal lines has been found by Lozitska & Lozitsky (1982, 1988). The whole flare emission could be easily divided into polarized and non-polarized components. The relation between the equivalent widths of such emission and the measured magnetic field indicated the existence of many subtelescopic flux-tubes in the flare, which rapidly expanded with height. Qualitatively, such a field structure agrees with that derived by Keller *et al.* (1990) using an inversion of Stokes profiles from a plage and a network region. A more complicated case was found by Lozitsky *et al.* (2000) in another flare. A two-component field structure with a non-monotonous height distribution of the field strength in the atmosphere was discovered. A field maximum of 4 kG was found in the upper chromosphere or the temperature minimum. Penn & Kuhn (1995) observed the chromospheric line HeI 10830 Å in three flare kernels and found an emission line profile and a mean magnetic field of 0.74 kG; the exact height of formation of this line is unknown.

In this paper, we make a comparative study of two solar flares using several methods which could give some information about the local magnetic field strength at the photospheric level. A first study of that type was presented at the Polarization Workshop 3 and published in a short paper by Lozitsky & Staude (2003).

2. Observational data and selected lines

The observational data were obtained with the Echelle spectrograph of the horizontal solar telescope of the Kyiv University Astronomical Observatory (Kurochka *et al.* 1980). This instrument can record the solar spectrum simultaneously from 3800 to 6600 Å with a spectral resolution of nearly 200,000 in the green region and with a temporal resolution of about several seconds.

Table 1. List of selected spectral lines.

Number	Wavelength (Å)	Element and multiplet number	Equivalent width W (mÅ) Rowland table	Excitation potential (eV)	Effective Lande factor g_{eff}	Height of formation (km)
1	5232.946	FeI-383	346	2.94	1.261	–
2	5247.052	FeI-1	59	0.09	1.998	328
3	5247.564	CrI-18	76	0.96	2.50	308
4	5250.212	FeI-1	62	0.12	2.999	324
5	5250.650	FeI-66	104	2.20	1.502	330
6	5576.097	FeI-686	113	3.42	–0.012	–

The flare of 25 July 1981 had the coordinates 11°S , 36°E and the importance $2N$. Three photographic Zeeman spectrograms were obtained with a circular polarization analyzer at 12:48, 12:58, and 13:25 UT. The first spectrogram refers to the phase near the flare peak and the last one to the end of the flare. All spectrograms refer to the same position on the Sun where the heliocentric angle was $\mu \approx 0.77$. For all spectrograms the entrance slit of the spectrograph was crossing the position of two bright flare knots outside the large spot in an area of magnetic fields with both S - and N -polarities. In the analysis presented below, we shall discuss the place with the strongest flare emission with a diameter of about 3000 km and magnetic N polarity.

The second flare of 5 November 2004 occurred in AR 10696 at $\mu \approx 1$ with an importance of M4.1/1B. Six Zeeman spectrograms were obtained between 11:35:35 and 11:49:00 UT. In the present study we analyze the spectra of 11:36:55 UT corresponding to the peak of the flare, considering a place in the tail part of the delta configuration, close to the sunspot of N polarity with a field strength of 2500 G in the umbra. All spectra were made with ORWO WP3 photo-emulsion.

Six spectral lines were used for the magnetic field diagnostics (Table 1). The lines, nos. 2–5 are well known, they are formed in the narrow range of middle photosphere – at a height of $h = 308\text{--}330$ km (Gurtovenko & Kostik 1989). This circumstance is important for small-scale magnetic field diagnostics (Stenflo 1973). In this connection, it is important and useful that these lines have Lande factors in a wide range from 1.5 to 3.0, with discrete step of about 0.5. In addition, their temperature sensitivities are also mutually close that follows from relatively close values of excitation potentials (0.09–2.20 eV).

For flare-active regions, values of formations can be somewhat different, but we do not expect strong shifts of the height of line formation because in the lower and middle photosphere, the thermodynamic conditions for flares of importance 1 and 2 are practically similar to those in the undisturbed atmosphere (Baranovsky *et al.* 1991; Lozitsky & Baranovsky 1993).

The spectral line FeI 5232.9 is much stronger than the other lines, its core is formed near the temperature minimum ($h \approx 500$ km), although the formation height of the whole line is close to that of the other lines, nos. 2–5. We add this line to check some fine spectral effects discovered first in the other magneto-sensitive lines. Analogously, we have investigated the FeI 5576.1 line to study the non-magnetic effects in the line profiles. Its level of formation is also close to lines nos. 2–5.

The effective Lande factors g_{eff} for all lines excluding line no. 3 have already been determined in the laboratory (Zemanek & Stefanov 1976; Landi Degl’Innocenti 1982),

which is important for reliable magnetic field measurements (Stenflo *et al.* 1984). The value of g_{eff} for line no. 3 in Table 1 corresponds to the theoretical case of LS coupling.

3. Magnetic field diagnostics

3.1 Averaged longitudinal magnetic field $B_{\parallel}$

The simplest test of the presence of unresolved magnetic fields consists of a comparison of the longitudinal magnetic fields $B_{\parallel}$ measured in several spectral lines with different Lande factors but formed practically at the same depths and having similar temperature sensitivities. Such magnetic field strengths $B_{\parallel}$ averaged over the aperture can be measured as the relative shift of the ‘centers of gravity’ of the $I + V$ and $I - V$ profiles. The physical meaning of this parameter is close to that of the longitudinal magnetic flux, when $\Delta\lambda_H/\Delta\lambda_D \ll 1$, where $\Delta\lambda_H$ is the Zeeman splitting and $\Delta\lambda_D$, the Doppler half width. If the magnetic flux is intermittent in the form of flux-tubes with a field strength B and a filling factor α , then:

$$B_{\parallel} \approx \alpha B \cos \gamma, \quad (1)$$

where γ is the angle of field inclination with respect to the line-of-sight.

If the flux-tube magnetic field is very strong, $\Delta\lambda_H/\Delta\lambda_D \geq 1$, then the observed field, B_{obs} is

$$B_{\text{obs}} < B_{\parallel} = \alpha B \cos \gamma, \quad (2)$$

due to the magnetic saturation effect (Howard & Stenflo 1972). Exactly, $B_{\text{obs}} < B_{\parallel}$ is real only in the case of measurements with solar magnetograph which have fixed position of the exit slits. On the contrary, if we measure using photographic method (as in the present work), we can obtain $B_{\text{obs}} = B$ due to the direct observations of fully-splitting Zeeman components.

Our spectral measurements in the lines, nos. 2–5 for flares are given in Fig. 1 together with similar data for weak magnetic fluxes, $B_{\parallel} < 100\text{--}200$ G (Gopasyuk *et al.* 1973) and with spectral data for moderate fields ($B_{\parallel} \approx 500$ G) in an active region outside flares (Lozitsky 2003). We can see very different dependences: in general, increasing relative fields $B(g_i)/B(g = 3.0)$ with decreasing Lande factors in first and second cases, but opposite tendency – decreasing relative values $B(g_i)/B(g = 3.0)$ with decreasing Lande factors – in the case of both flares. Note that measured magnetic fluxes in both flares were also moderate, in the range of 250–700 G.

Beginning the interpretation of results presented in Fig. 1, let us put the question: which shape of dependence “ $B_{\parallel}$ versus g ” must be in a case of weak or moderate (<500 G) homogeneous field?

In this case, Stokes V parameter could be presented *via* a formula

$$V \propto \left(\frac{dI}{d\lambda} \right) \Delta\lambda_H, \quad (3)$$

where $dI/d\lambda$ is the Stokes I gradient and $\Delta\lambda_H$ is Zeeman splitting.

$$\Delta\lambda_H = 4.67 \times 10^{-13} g \lambda^2 B, \quad (4)$$

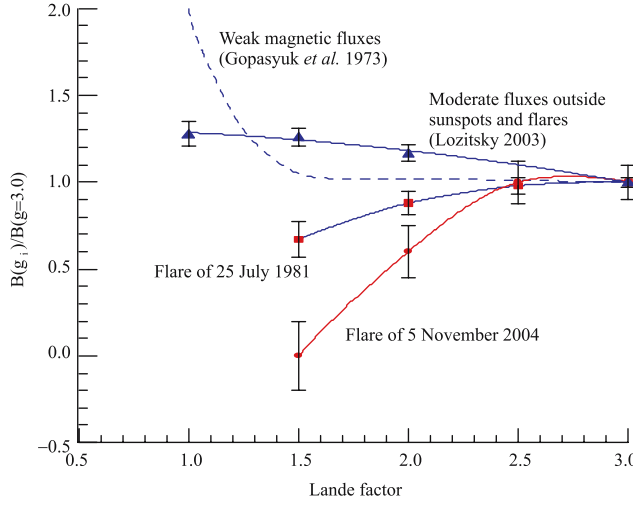


Figure 1. Comparison of the relative observed magnetic field values $B(g_i)/B(g = 3.0)$ for weak magnetic fluxes ($B_{\parallel} < 100\text{--}200$ G) from magnetographic measurements (Gopasyuk *et al.* 1973), for moderate fluxes ($B_{\parallel} \approx 500$ G) measured in active regions outside flares (Lozitsky 2003), and for studied flares.

where $\Delta\lambda_H$ and λ are in Å, and B in gauss (G). In practice, measured magnetic field strength $B_{\parallel}$ can be found as:

$$B_{\parallel} \propto \frac{V}{g}. \quad (5)$$

From equation (3)–(5) it follows that *in case of weak or moderate field, values of $B_{\parallel}$ do not depend on Lande factor g .*

Second possible situation is the following: two-component field contains weak background field and small-scale moderate field. In this case, observed Stokes V_{obs} picture forms as a result of interflowing two Stokes V ‘waves’: first from background field V_{backgr} and the second from small scale (perhaps flux-tube) field V_{fluxtube}

$$V_{\text{obs}} = V_{\text{backgr}} + V_{\text{fluxtube}}. \quad (6)$$

Both these ‘waves’ have very similar spectral distributions, but different amplitudes. Amplitude of each ‘wave’ is proportional to filling factor of each magnetic component, and also Lande factor g of line. If we use different spectral lines with different Lande factors, measured value of $B_{\parallel}$ should be, as first approximation, *the same for different lines, independently from Lande factor*. In this case we can expect $B(g_i)/B(g = 3.0) \approx 1$ in Fig. 1.

The third possible situation: two-component structure, with weak or moderate background field and small-scale (spatially unresolved) strong (>1 kG) field. In this case, position of Stokes V peaks of background field, V_{backgr} , is practically the same for lines with different Lande factors. It corresponds, in general, to peaks of Stokes I gradient distribution, $dI/d\lambda$. On the contrary, position of Stokes V peaks of strong small-scale field, $V_{\text{flux-tube}}$, must be very different and should be close to the Zeeman splitting $\Delta\lambda_H$ in the second component. For spectral lines with high Lande factors, e.g.,

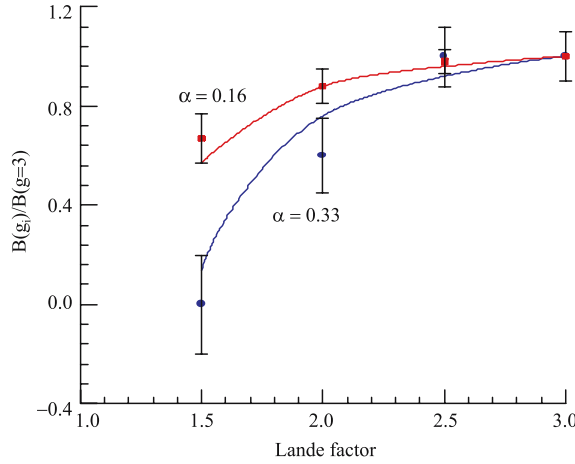


Figure 2. Comparison of the observational and theoretical values of $B(g_i)/B(g = 3.0)$ versus Lande factor for both flares. Squares and filled circles present observations for flares of 25 July 1981 and 5 November 2004, respectively; solid curves – theory for different filling factor (see text).

$g = 2.5$ – 3.0 , two named Stokes V waves are mutually ‘splitted’, i.e., non-blended. If we use another line with some smaller Lande factor, for instance, $g = 2.0$, the Stokes $V_{\text{flux-tube}}$ wave begins to blend far wings of first Stokes V wave, and it leads to change of measured magnetic field strength $B_{\parallel}$. If magnetic polarities in both components are the same, we should obtain $B_{\parallel}(g = 2) > B_{\parallel}(g = 2.5$ – $3)$. On the contrary, if these polarities are opposite, we obtain $B_{\parallel}(g = 2) < B_{\parallel}(g = 2.5$ – $3)$ due to mutual cancellation of circular polarization of opposite signs. We can expect analogous results also in case, if these polarities are the same, but Zeeman manifestations are different; absorption type for one component and emission for second one. Also, we can expect the maximum effect of $B_{\parallel}$ increasing (or decreasing) in the case when both above-mentioned Stokes V waves interflow fully, namely, in lines with very small Lande factors.

Figure 2 presents a comparison of our observations with calculations. A simple two-component magnetic model was assumed, with a field strength of $B_{\text{backgr}} = +500$ G in the first background component, and $B_{\text{fluxtube}} = -4000$ G in the second small-scale one. The filling factors of these components are α_1 and α_2 , respectively, where $\alpha_1 + \alpha_2 = 1$. The Stokes I and V profiles were calculated for a Milne-Eddington atmosphere model using Unno’s (1956) parameters η_0 and $\Delta\lambda_D$ close to those of a line such as FeI 5250.2, that is 2.5 and 40 mÅ, respectively. It was also assumed that in the area of the strong field component the Doppler half width of the line profiles is about 30% smaller than that in the first background field component. A similar effect of narrow line profiles has been observed earlier, e.g., by Lozitsky (1980) and Lozitzka & Lozitsky (1982). Magneto-optical effects were neglected because these effects are important in the central parts of the profiles only, i.e., for distances of less than 40 mÅ from the line centers (Lozitsky & Sheminova 1992). However, the most interesting effects for our study occur in the far line wings.

We can see in Fig. 2 that a filling factor of $\alpha_2 \approx 0.16$ is fitting the first flare, and $\alpha_2 \approx 0.33$ to the second one. Thus, the calculations give indications of roughly the same

magnetic field strength of $B_2 \approx -4000$ G in the subtelescopic structures of both flares, but essentially different (about 2 times) filling factors. Note that the accuracy of our simple calculations is about 5–10%, and thus we neglect possible fine differences for these flares.

So, we can conclude that the described simple test based on $B_{\parallel}(\lambda_i)$ values for different Lande factors gives an observational evidence to two possibilities:

- possible presence of strong unresolved magnetic fields of opposite polarity and
- presence of strong unresolved fields of same polarity, but in spectral manifestation of the emissive Zeeman effect.

Note that the last case was not visible directly: all investigated Stokes $I \pm V$ profiles (for all lines listed in Table 1) were smooth, without emission peaks in their cores. If possible emission features in profiles are really existing, ones should have small intensities due to likely small filling factor.

3.2 The 5250.2/5247.1 line ratio

The magneto-sensitive lines FeI 5250.2 and 5247.1 have practically the same heights of formation and of temperature sensitivity, which is especially important for small-scale magnetic field diagnostics. If subtelescopic magnetic kG fields really exist, we can expect non-constant line ratio $B_{\parallel}(5250.2)/B_{\parallel}(5247.1)$ at different distances, $\Delta\lambda$, from line center (Stenflo 1973). Because these lines have practically the same Stokes I profiles, instead of ratio $B_{\parallel}(5250.2)/B_{\parallel}(5247.1)$ we can consider as equivalent Stokes V ratio $2V_{5250.2}/3V_{5247.1}$ (Stenflo *et al.* 1987). As to theory, we can expect $2V_{5250.2}/3V_{5247.1} \approx 1$ in case of weak or moderate fields inside the aperture. On the contrary, if this ratio is far from unity, the magnetic fields should be strong.

Figure 3 displays our results for the flares in comparison with the analogous data for solar faculae by Frazier and Stenflo (1978). We can see surprisingly good agreement in both data for faculae and flare of 25 July 1981, but essential difference named data *versus* data for flare of 5 November 2004 (typical error ± 0.05).

So long as our results for flare of 25 July 1981 are very similar to the one for faculae, we can use the theoretical model by Frazier & Stenflo (1978) for the interpretation of our data. Frazier & Stenflo concluded that as pointed out in Fig. 2, dependence for faculae could indicate the presence of small-scale magnetic fields of 2.0 kG strength assuming non-rectangular field profile. In the case of rectangular profile magnetic field, strengths should be less in 1.5–2 times (Stenflo 1973; Lozitsky & Thap 1989).

As to flare of 5 November 2004, we can expect also kG fields in small-scale features, as it follows from comparison of these observations with theoretical dependences published in a paper by Rachkovsky *et al.* (2005). However, true amplitude of small-scale magnetic field should be here on about 300–500 G more, as at first flare.

The third test of small-scale magnetic fields is based on FeI 5247.1 and 5250.2 lines too. The particular drawback of these lines is their relatively high temperature sensitivity (Harvey & Livingston 1969; Staude 1970a, b). Lozitska & Lozitsky (1994) have shown that the temperature sensitivity of the pair 5250.2/5247.1 is not an obstacle for reliable magnetic field measurements if we do not use the Stokes V amplitude (such as in a solar magnetograph), but the Stokes V peak separation, $\Delta\lambda_V$. Also, the Stokes V peak half-width, $\Delta\lambda_{1/2,V}$ and the Stokes I half-width, $\Delta\lambda_{1/2,I}$, could be successfully used for small-scale magnetic field diagnostics.

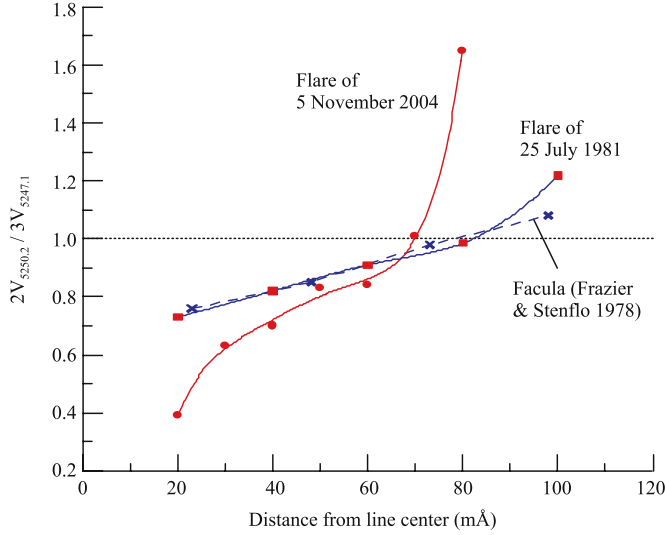


Figure 3. Comparison of Stokes V ratio $2V_{5250.2}/3V_{5247.1}$ in FeI 5250.2 and FeI 5247.1 lines for the flares (present study) and faculae (Frazier & Stenflo 1978). We can see that, in general, $2V_{5250.2}/3V_{5247.1} \neq 1$ and this circumstance indicates the presence of small-scale kG fields in solar flares.

Table 2. Observed Stokes I half-width, Stokes V peak separation, and Stokes V peak half-width.

Line	$\Delta\lambda_{1/2,I}$	$\Delta\lambda_V$	$\Delta\lambda_{1/2,V}$
FeI 5247.052	133	106	79
FeI 5250.212	161	126	80

The observed values of the mentioned parameters (in mÅ) for the flare of 25 July 1981 are given in Table 2.

Let us consider a simple model according to which small-scale flux-tubes with a field strength, B_f are embedded in a background field, B_b . The calculation scheme for such a model was described in detail by Lozitska & Lozitsky (1994) and Lozitsky *et al.* (2000).

If we assume a relative absorption coefficient, η_0 , close to 2.5 and an angle of inclination with respect to the line of sight, γ , between 0° and 90° , then from the ratio

$$\frac{\Delta\lambda_{1/2,I}(5250.2)}{\Delta\lambda_{1/2,I}(5247.1)} = 1.21, \quad (7)$$

follows $\Delta\lambda_D = 50$ mÅ, $\Delta\lambda_H = 40$ mÅ for $\gamma = 0^\circ$, and $\Delta\lambda_D = 48$ mÅ, $\Delta\lambda_H = 50$ mÅ for $\gamma = 90^\circ$. The first case corresponds to $B_b = \pm 1.05$ kG, and the second one to $B_b = 1.3$ kG. The first case means that we observe a dispersion of field directions, with spatially unresolved contacts of background magnetic elements of both (N and S) polarities. We can compare the derived values with the typical dispersion of about ± 0.5 kG for non-spot and non-flare regions (Harvey *et al.* 1972), and ± 0.9 kG for a 2B-flare (Lozitsky *et al.* 2000).

Taking into account the Stokes V peak ratio

$$\frac{\Delta\lambda_V(5250.2)}{\Delta\lambda_V(5247.1)} = 1.19, \quad (8)$$

we have $B_f = 1.45$ kG, if $\gamma = 45^\circ$, i.e., close to the heliocentric distance for the studied flare. It is interesting to note that the derived value of $B_f = 1.45$ kG is close to the value $B_1 \approx 1.3$ kG obtained from the $(BS - \Delta\lambda/g\lambda^2)$ diagram (see sections 3.5 and 3.6).

However, the Stokes V peak half-widths, $\Delta\lambda_{1/2,V}$, could not be explained on the basis of such a two-component model. Theoretically, we have $\Delta\lambda_{1/2,V} = 76.5$ mÅ for FeI 5250.2, but 80 mÅ from observations. For the FeI 5247.1 line, the theoretical value is 67 mÅ, whereas the observed one is 79 mÅ. This difference (12 mÅ) is too large and needs some modification of the present model. One possible solution is an additional third magnetic field component which broadens the Stokes V peak. In fact, if the Zeeman splitting $\Delta\lambda_H$ in such a third component is close to 150 mÅ (i.e., $B \approx 3.9$ kG for FeI 5250.2), then for FeI 5247.1 the same field strength products $\Delta\lambda_H \approx 100$ mÅ due to the Lande factor ratio of 1.5. So long as $\Delta\lambda_H \approx 100$ mÅ corresponds to the middle level of the Stokes V peak for lines like FeI 5250.2, we can expect an essential influence on the peak extension, in this case. On the contrary, the value of $\Delta\lambda_H \approx 150$ mÅ corresponds to the Stokes V wing (no middle level), that would not cause an increase of the $\Delta\lambda_{1/2,V}$ value.

Thus, the Stokes V peak half-width data give evidence of the existence of *more than two* structure components in the flare magnetic field.

3.3 Stokes V profiles

The fourth test of small-scale fields is based on the study of Stokes V profiles too. As it was pointed above (see equation 3), for really weak magnetic field ($\Delta\lambda_H/\Delta\lambda_D \ll 1$) we can expect that $V \propto (dI/d\lambda)\Delta\lambda_H$. From this equation, it follows that the spectral distributions of the Stokes V parameter and the Stokes I gradient $dI/d\lambda$ coincide, if the magnetic field is really weak (less than 1 kG).

Some of the observed cases are given in Figs. 4 and 5. We can see some deviations between V and $dI/d\lambda$ which have the tendency to be placed almost symmetrical to the line centers. Of course, this could be by chance, and we must check the noise level to obtain a more reliable conclusion.

To estimate these effects, we made photometry of spectral continuum in the range of 5248.5–5248.8 Å and found that typical intensity fluctuations connected with grain of photo-emulsion is relatively weak, average of 1%. So, we can expect that differences between V and $dI/d\lambda$ exceeding 1% could be real.

On the other hand, some details presented in Figs. 3 and 4 cover a spectral range of about 30–50 mÅ, i.e., they are close to instrumental fluctuations. In principle, this does not exclude the reality of the effects under discussion. Lozitsky (1980) discovered similar narrow spectral features in an active region outside the flares using many different magneto-sensitive lines. Lozitsky *et al.* (1999) observed in a solar flare relatively weak (5–10%) emission peaks in FeI line cores which had observed half-widths of about 30–40 mÅ, but after reduction for instrumental broadening, it was about 10–20 mÅ only! Obviously, the lower limit of the half-widths of spectral lines in volumes with

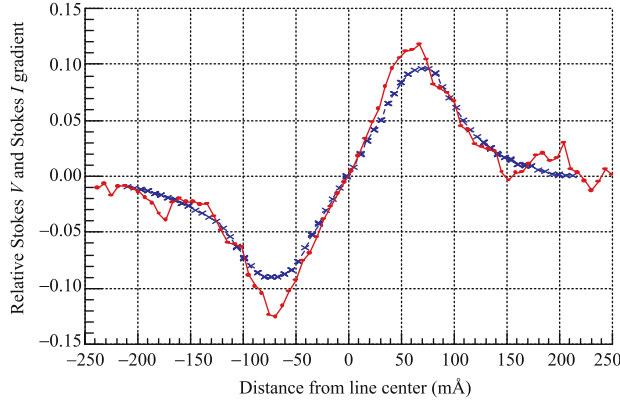


Figure 4. Observed Stokes V profile (solid line) and Stokes I gradient $dI/d\lambda$ (crosses and dashed line) for the FeI 5250.6 line in the flare. One can see the red–blue asymmetry of Stokes V and some differences between the parameters, in particular at $\Delta\lambda = \pm 70$ and at $\pm (180\text{--}200)$ mÅ.

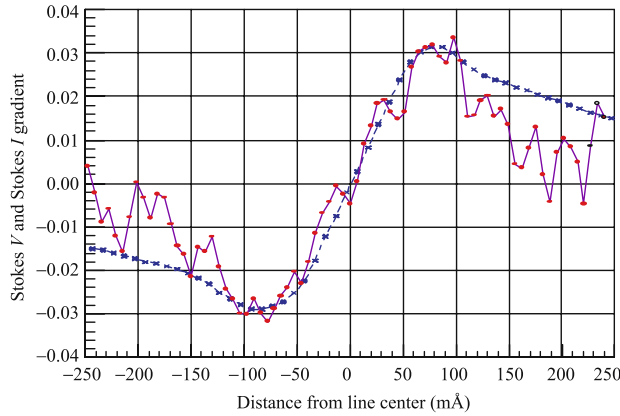


Figure 5. Observed Stokes V profile (solid line) and Stokes I gradient (crosses and dashed line) for FeI 5233 line in the flare of 25 July 1981.

very strong and small-scale fields is practically unknown at present, and therefore we could continue the analysis for obtaining more clear conclusions.

3.4 Bisector splitting

For the flares and spectrograms under study, there is only one possible way to intensify the suspected magnetic effects in the line profiles. We could investigate a spectral line with a larger equivalent width W . In this case, the possible ‘magnetic’ spectral features should be more intensive, whereas the non-magnetic fluctuations are unchanged. A suitable line is FeI 5233, which has $W = 346$ mÅ.

In addition, it is interesting to study the ratio

$$BS = \frac{V}{dI/d\lambda}. \quad (9)$$

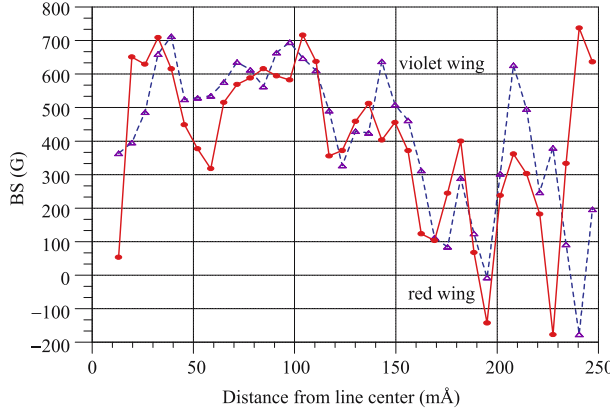


Figure 6. Comparison of bisector splitting BS for violet (dashed) and red (solid) wings of FeI 5232.9 line for the flare of 25 July 1981. We can see the close correlation between named values, especially for distances less than 230 mÅ. The correlation coefficient r is 0.72, and the probability p of an accidental connection is 0.001.

This parameter BS , the ‘bisector splitting’, is similar to the Zeeman splitting $\Delta\lambda_H$ for any given distance from the line center $\Delta\lambda$, see equation (3).

Figure 6 displays an example of the observed $BS - \Delta\lambda$ distribution for the flare of 25 July 1981. We can see strong fluctuations of BS which display the tendency to be disposed by pairs with respect to the line center. In particular, there are pairs of maxima at $\Delta\lambda = \pm (30-40)$, ± 100 and perhaps ± 140 mÅ. The minima of BS are disposed at $\Delta\lambda = \pm 55$ and at $\pm (120-130)$ mÅ.

Analogous effects were found in other lines, and in the flare of 5 November 2004 too. It is important to note that such effects are impossible in the case of weak or moderate field. If the magnetic field strength really equals 700 G, we can expect

$$BS(\Delta\lambda) \approx \text{const.} \quad (10)$$

for all investigated lines. Of course, in this case there should be no pairs of spectral features, and all fluctuations should be irregular.

3.5 $(BS - \Delta\lambda/g\lambda^2)$ diagram

So long as the values of BS for the red and violet wings are mutually correlated, we can calculate the average values of BS for both wings of each line. In addition, if each statistically reliable maximum in the $(BS - \Delta\lambda)$ diagram disposed at a distance of $\Delta\lambda_{\max}$ indicates a Zeeman σ -component related to a small-scale magnetic feature, then

$$\Delta\lambda_{\max} = \Delta\lambda_H = 4.67 \times 10^{-13} g\lambda^2 B, \quad (11)$$

where $\Delta\lambda_H$ and λ are in mÅ, and B in gauss (G).

From equation (11) it follows for $B = \text{const.}$, that

$$\frac{\Delta\lambda_{\max}}{g\lambda^2} = \text{const.} \quad (12)$$

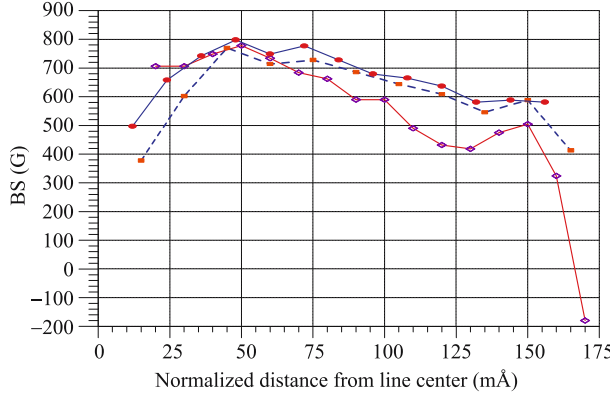


Figure 7. Bisector splitting BS versus the value $82.7\Delta\lambda/g\lambda^2$ for three lines: FeI 5250.2 (rhombs and solid line), FeI 5247.1 (dashed line), and CrI 5247.6 (filled circles and solid line).

The criterion (12) can be used additionally to verify the magnetic nature of the spectral peculiarities (Lozitsky 1980).

A typical example of the observed $(BS - \Delta\lambda/g\lambda^2)$ diagram for flare of 25 July 1981 is shown in Fig. 7. All BS values were calculated for normalized distances from the line center, $\Delta\lambda_n$, according to

$$\Delta\lambda_n = \frac{82.7\Delta\lambda}{g\lambda^2}. \quad (13)$$

So long as the factor $g\lambda^2 = 82.7 \times 10^6$ corresponds to the FeI 5250.2 line, all data were reduced to the case of that line.

We can see a good agreement for the three lines, nos. 2, 3, and 4 with the largest Lande factors (Fig. 6). The first common maximum corresponds to $\Delta\lambda_{\max}(1) \approx 50 \text{ mÅ}$, and the second one to $\Delta\lambda_{\max}(2) \approx 150 \text{ mÅ}$. The FeI 5250.6 line with Lande factor $g = 1.5$ has also a similar distribution of bisector splitting, with deep minima on 300 mÅ of normalized distance from line center. This minima, reduced to the case of FeI 5233 line and non-normalized distance, corresponds to

$$\Delta\lambda_{\min}(5233) = 300 \times \left(\frac{1.261}{2.999} \right) \approx 126 (\text{mÅ}). \quad (14)$$

(Note that here values of 1.261 and 2.999 are Lande factors for FeI 5233 and FeI 5250.2 lines.) As to reality, one can see from Fig. 6, that FeI 5233 line, in fact, has BS minimum on close distance $\approx 120 \text{ mÅ}$. Thus, we can conclude that BS distributions for different lines are in satisfactory agreement, and this circumstance gives the evidence to the magnetic nature of common extremums in Fig. 7.

An additional argument to this point of view follows from the comparison of data for both flares (Fig. 8). We can see that two investigated flares give very similar and strong BS distributions, which are obviously greater than possible non-magnetic effects observed in ‘non-split’ FeI 5576 line. In the last case, it was formally supposed that its Lande factor is also 1.261, as in FeI 5233 line (really, it is -0.012 , i.e., in ≈ 100 times less). One can see the close correlation between BS distributions for both

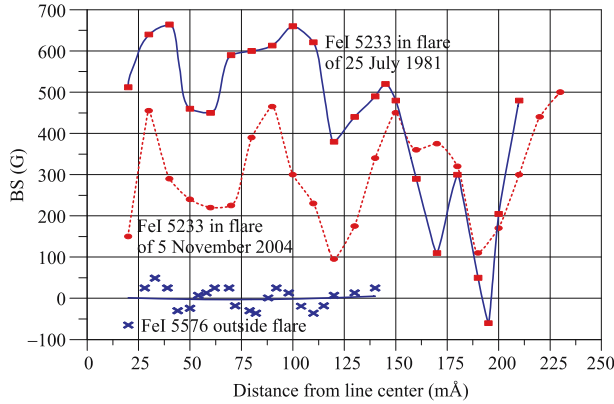


Figure 8. Comparison of bisector splitting in FeI 5233 for flares of 25 July 1981 (solid) and 5 November 2004 (dashed). For estimation of the possible instrumental effects, the equivalent BS fluctuations in ‘non-split’ FeI 5576 line (crosses) are presented too (see text).

flares, and common minima on the following distances $\Delta\lambda = 60, 120$ and perhaps 190 mÅ .

Comparison of FeI 5233 and FeI 5576 indicates that at least two first minima on distances $\Delta\lambda = 60, 120 \text{ mÅ}$ should be real and have no magnetic nature. In fact, possible standard deviation of measurements should be on a level about 50 G , as it follows from FeI 5576 line. But observed BS fluctuations in FeI 5233 reach out of $200\text{--}300 \text{ G}$, i.e., they are much stronger. So, we do not find any non-magnetic cause which could give similar perturbations.

3.6 Possible magnetic field strengths

In general, to obtain the true magnetic field strengths in unresolved magnetic structures, the magnetic polarity and type of the Zeeman effect, namely emission and absorption, should be made out. For each diagnostic diagram as shown in Fig. 7, we cannot conclude which type of extremum (minimum or maximum?) corresponds to the position of contribution of the Zeeman sigma-components in the spectra.

On the contrary, we can attempt a more reliable conclusion about the presence of *several* types of magnetic features having *different* magnetic field strengths. Also, we can point possible magnetic field range observing the spectral range of normalized distance from line center (see Fig. 7) where the correlated fluctuations of BS parameter were measured.

In principle, the determination of the magnetic polarity and type of Zeeman effect in such a complicated case could be determined from the comparison of Stokes V and Stokes I distribution. In particular, if we really observe the Zeeman effect in absorption for the same (N) polarity, we should discover very weak depression in Stokes I for the same $\Delta\lambda$ position where the BS maximums were found. On the other hand, if a very weak emission in Stokes I presents where the BS minimum exists, we have the emissive Zeeman effect for the same magnetic polarity.

For the 25 July 1981 flare, Lozitsky & Staude (2003) found a direct evidence of such a situation. The difference between the Stokes I profiles of the FeI 5250.2 line in the flare and at a place outside indicated the presence of *emissive* Zeeman sigma

components, while *the same* (N) polarity existed at the mentioned spectral position. This spectral position corresponds to a magnetic field strength of 7.8 kG.

An application of criterion (11) to the above listed values of $\Delta\lambda_{\max} = 50 \text{ mÅ}$ and $\Delta\lambda_{\max} = 150 \text{ mÅ}$ (Fig. 7) gives the following magnetic field strengths: $B_1 \approx 1.3 \text{ kG}$, $B_2 \approx 3.9 \text{ kG}$. Note that, second value is very close to the $B_{\text{true}} \approx 4.0 \text{ kG}$ which follows from Fig. 2 (see section 3.1).

Let us put the question: which spectral peculiarities should correspond to this magnetic field strength (3.9–4.0 kG) in Fig. 8 for FeI 5233 line? First maximum on $\Delta\lambda_{\max} \approx 35 \text{ mÅ}$ corresponds to 2.15 kG, but first minimum (on $\Delta\lambda_{\min} \approx 60 \text{ mÅ}$) corresponds to $B \approx 3.7 \text{ kG}$. Likely, minimum on $\Delta\lambda_{\max} \approx 60 \text{ mÅ}$ has a real physical meaning. In any case, results of three different methods are very close, if we consider that namely, minimum on $\Delta\lambda_{\min} \approx 60 \text{ mÅ}$ (Fig. 6) indicates the position of very weak Zeeman σ -components from any subtelescopic structures with very strong field but small filling factor.

Third field value, B_3 , could be close to 7.8 kG, as it follows from FeI 5250.6 line (see above). It is interesting to note that practically the same value, 7.4 kG corresponds to second minimum on Fig. 8, $\Delta\lambda_{\min} \approx 120 \text{ mÅ}$. Relatively small difference between these values (7.4 and 7.8 kG) could be due to the errors of measurements or vertical gradient of the magnetic field. Analogously to the second value $B_2 \approx 4.0 \text{ kG}$, we can suppose that magnetic elements with $B_3 \approx 7.4\text{--}7.8 \text{ kG}$ also give the very weak spectral contribution in the form of *emissive Zeeman effect* for the same (N) magnetic polarity.

Three comments should be given to these values:

- Three different magnetic field values B_1 , B_2 , and B_3 mean a multi-component magnetic field structure, probably in the form of three types of magnetic features with different magnetic field strengths (Lozitsky 1980; Lozitsky & Staude 2003).
- Doppler and turbulent velocities in such structures should be close to zero, that follows from the negligible relative shift of the extrema in the violet and red wings (Fig. 6) and from very narrow (about 30–50 mÅ) spectral peculiarities.
- If such ‘super-strong’ magnetic fields (close to $\approx 10 \text{ kG}$) really exist, we can observe very weak manifestation of the Zeeman effect *even in ‘non-split’ line* FeI 5576 because its Lande factor, strictly speaking, is non-zero, -0.012 , i.e., about 105 times less than in FeI 5233. For instance, magnetic field of 10 kG should produce in FeI 5576 such Zeeman splitting, as magnetic field of about 100 G in FeI 5233.

Wonderfully, observations confirm this assumption (Fig. 9). We can see that weak splitting in FeI 5576 is observed really in the central part of line profile ($\Delta\lambda < 60\text{--}80 \text{ mÅ}$), and its calibrated value, in fact, is close to 100–150 G. Note that analogous effect was discovered earlier also in other flares (Lozitsky 1993, 1998).

4. Discussion and conclusion

Comparison of all the data derived above leads to the conclusion that a four-component magnetic field structure is possible, with field strengths of about $\pm 1.05 \text{ kG}$ in the background field, and of 1.3–1.45, ≈ 4.0 , and 7.4–7.8 kG (at $h \approx 300 \text{ km}$) in the small-scale, spatially unresolved magnetic elements.

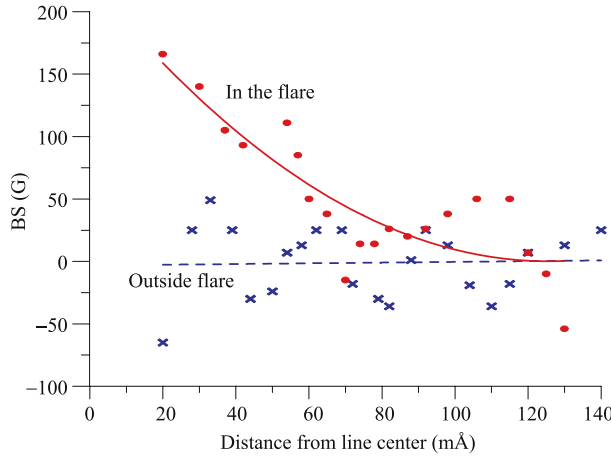


Figure 9. Observational manifestation of the Zeeman effect in core of FeI 5576 line in flare of 5 November 2004: crosses and dashed line – splitting of bisectors of $I \pm V$ profiles in this line, formally calibrated to the case of FeI 5233 line ($g_{\text{eff}} = 1.261$), outside the flare; filled circles and solid line – observed BS splitting in bright knot of flare.

As to the fields in the 1.1–2.3 kG range, they were measured earlier by many authors, e.g., by Stenflo (1973); Wiehr (1978); Koutchmy & Stellmacher (1978); Lozitsky & Tsap (1989) and Lozitsky *et al.* (2000). On the contrary, the flux-tube mode of ≈ 4.0 kG strength is almost unknown and unstudied. Lozitsky (1980, 1986) found observational evidence of this mode in non-flare regions. Obviously similar data were also obtained by other authors, but they probably considered the data as insufficient or unreliable.

Arguments for the possible existence of a magnetic mode with $B_3 \approx 8\text{--}10$ kG were given by Lozitsky (1979, 1980) for both flare and non-flare regions. At present we have only first preliminary data which could give hints at such ‘super-strong’ fields. In particular, it is unknown which is the upper field strength limit for the magnetic features. Observational evidences of flare fields of even 20–90 kG were found on the basis of a study of Stokes $I \pm V$ and $I \pm Q$ profiles in FeI, MgI and HI lines in a 2B-flare (Lozitsky 1993, 1998).

Why are these fields practically unknown?

Here are the main factors which make the diagnostics of such super-strong fields very difficult (see, e.g., Lozitsky 1993):

- A drop of the gas pressure in magnetic elements hampers their diagnostics, since it diminishes the optical thickness in the tubes. In the simplest case, when the flux-tube is homogeneous and untwisted, the condition

$$p + \left(\frac{B^2}{8\pi} \right) = p_{\text{ex}}, \quad (15)$$

should be met, where p and p_{ex} are the gas pressures inside and outside the tube, respectively; B is the magnetic field strength in the tube (the magnetic field pressure outside the tube is neglected). Clearly, the stronger B is inside the tube, the lower must p be when p_{ex} is fixed. We may therefore expect that magnetic

flux-tubes with the strongest fields will be most rarefied, i.e., *spectrally invisible*. Moreover, a magnetic field of 8 kG would require a $p_{\text{ex}} > 2.5 \times 10^6 \text{ dyn cm}^{-2}$; that is, 10^3 times larger than p at line-forming heights of the quiet Sun. The time scale of a pressure adjustment of such a huge local increase of p (if it exists due to the flare) would be *extremely short* and make the observation very difficult.

- A small value of the filling factor α : when $\alpha \ll 1$, the corresponding spectral features may be so feeble that it is practically impossible to separate them from the instrumental noise background or the emulsion grains. In this case, we must check carefully the spectral manifestation using special statistical methods.
- The dispersion of magnetic field strengths inside the flux-tubes can also mask the presence of super-strong fields. It should manifest itself in a smearing (diminishing depth and contrast) of the Zeeman σ -components.
- The diagnostics of super-strong magnetic fields can be made either difficult or easy depending on the local thermodynamic conditions. It is likely that the corresponding magnetic flux-tubes in flares are more conspicuous due to the increased gas pressure in them and to emission reversal of the Zeeman pattern. Flares might be the areas where super-strong fields are most likely to arise – there is empirical evidence that a local (vertically) intensification of the magnetic field (a ‘collapse’) occurs in flare areas (Lozitsky *et al.* 2000; Kurochka & Lozitsky 2005). Thus there is good reason to look for extra manifestations of super-strong fields just in flares. The diagnostics of the complicated magnetic structure in a flare could probably be improved by applying a Stokes profiles inversion procedure to the data (see, e.g., Carroll & Staude 2001).

Another interesting effect discovered two decades ago from observations (Lozitsky 1980, 2003) is a field strength ‘quantization’. It was shown that a theoretical interpretation of these phenomena can be offered within the frame of a linear force-free model (Soloviev & Lozitsky 1986). This model has a multi-polar periphery and a very strong field with discrete values near the tube axis. From a qualitative point of view, such a multi-polar periphery could give a spectral signature similar to a strong ‘mixed polarity’ background field. The theoretical values of discrete strengths near the tube axis are in satisfactory agreement with the observations.

It seems very much like that idea about possible magnetic field strength discretization in small-scale magnetic elements begins to be in the air. Recently Socas-Navarro & Lites (2004) found evidences for mixed field strengths in the quiet Sun, with two discrete values at 300 and 1700 G. Socas-Navarro (2004) obtained several magnetic strength ‘nodes’ in small-scale magnetic elements. He applied multi-line Stokes analysis using observational material from three modern spectropolarimeters (THEMIS, SPINOR and TIP + POLIS). Perhaps, some general peculiarities of the small-scale magnetic fields (including magnetic field strength discretization) are similar for flare and non-flare places on the Sun.

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