Solar Chromospheric Flares: Observations in Ly-α and Hα and Radiative Hydrodynamic Simulations

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The most exciting phrase to hear in science, the one that heralds new discoveries, is not “Eureka!” but “That’s funny...”

— Isaac Asimov
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Abstract

This thesis is divided into two main parts: a multi-wavelength observational study of solar flares, focusing mainly in the chromosphere in Ly\(\alpha\) and H\(\alpha\), and an application of a radiative transfer code and a radiative hydrodynamic code, to compare the results obtained by observations with the simulated ones.

The Ly\(\alpha\) emission is a very interesting line because it is a natural tracer of the solar activity in the chromosphere. The Transition Region And Coronal Explorer satellite observed a small number of flares in the Ly\(\alpha\) passband, but apart from this, these events have not often been observed in this strong chromospheric line. Because TRACE has a broad Ly-\(\alpha\) channel, in order to estimate the “pure” Ly\(\alpha\) emission, we had to apply an empirical correction.

We found that there is a reasonable coverage in TRACE 1216 \(\text{Å}\) and the TRACE 1600 \(\text{Å}\) for two different flares: on 8 September 1999 and on 28 February 1999. Studying them we estimated, for the first time, the pure Ly\(\alpha\) flare signature, being on the order of \(10^{25} \text{ erg s}^{-1}\) at the flare peak.

The study of the first flare gave us the possibility to calculate the electron energy budget using the X-ray data from Yohkoh/HXT in the context of the collisional thick target model, finding that the Ly-\(\alpha\) power is less than 10% of the power inferred by the electrons.

The morphology and evolution of the second flare were described in different wavelengths by using imaging data acquired by TRACE and by BBSO in white light and in H\(\alpha\). We studied the magnetic topology using the magnetic field provided by SOHO/MDI, extrapolating the photospheric magnetic field lines, assuming a potential field. We found different morphologies in the magnetic configuration before and after the flare, confirming the occurrence of a reconnection process.

The H\(\alpha\) line is the most important line in the chromosphere. We studied the H\(\alpha\) emission of a flare which occurred on 3 July 2002 using some spectroscopical observations from the Ondřejov Observatory. Analyzing the available data in other wavelengths, we made a morphological study of the active region from three hours before the flare to seven hours after it.

The results obtained by observations, both in the form of integrated intensity as a function of time, and detailed line profiles, motivated the second part of the thesis. In this, we used a radiative transfer code \cite{Gouttebroze:1978} applying different atmospheric models as input parameters in order to compute the hydrogen spectral lines and study...
how they are affected by the temperature and microturbulent stratification. In particular, the intensity of the Lyα and Hα lines is related to the temperature stratification of the atmospheric model, the position of the transition region being a key factor. The variation of the microturbulent velocity does not significantly affect the resulting intensities, but we observed that an increase of the microturbulent velocity broadens the line profiles.

The RADYN Radiative HydroDynamic code (Allred et al. 2005) was applied to solar flares, modelling a flare loop from its footpoints in the photosphere to its apex in the corona by adding non-thermal heating at the lower atmosphere and soft X-ray irradiation. The majority of this work was to deal with investigating the dynamical response of the solar chromosphere to energy injected in the form of non-thermal electrons during solar flares. We studied the flare energy transport and radiation production in the chromosphere as well as the Hα and Lyα emission. The Lyα intensity is affected by the flux of the initial beam of electrons injected at the top of the loop, while the Hα intensity appears to be less affected by the flare model.

Comparing the observational results in Lyα and Hα with the computed ones from the radiative code and the RADYN code, we found that the RADYN code fits better the Hα intensities to the observations than the Lyα intensities, concluding that the code gives a better description of processes in the lower chromosphere than those in the upper layers.
INTRODUCTION
Chapter 1

Introduction

The Sun is our nearest star: this fact allows us to study it in detail, discovering a vast number of interesting physical phenomena.

During the last years, solar physicists have increased the knowledge of both the solar interior and the solar atmosphere, recognizing that solar phenomena are much more complex than we could imagine.

The Sun is a unique laboratory for understanding fundamental physical processes such as particle acceleration, plasma heating, magnetic instabilities, reconnection and turbulence. We can thus consider the Sun as the Rosetta Stone of astrophysics. Methods originally developed for the Sun, as for example modelling of the interior and of the atmosphere, are now applied to other stars.

The existence of the Sun has a considerable influence on the life on Earth, i.e. high energetic events occurring in the solar atmosphere like coronal mass ejections, which can influence the performance and reliability of space-borne and ground-based technological systems. All these variety of processes which influence in our space environment are known as Space Weather. An understanding of space environmental conditions is also important in designing shielding and life support systems for manned spacecrafts.

Solving the problems relevant to solar physics requires a large coordinated effort in connecting together many different strands such as multi-wavelength observations, theory and high performance computing.

1.1 Thesis Summary

This thesis deals with the study of solar flares in the chromosphere and is divided into three main parts: the introduction (Chapters 1, 2 and 3), observations of different solar flares (Chapters 4 and 5) and simulations using two different codes (Sections 6 and 7).

In Chapter 1 there is a short generic introduction about the Sun and its structure.

Chapter 2 reviews the necessary background relevant to this thesis, from the description of its morphology and evolution to the different models.

In Chapter 3 we describe the characteristics of the instrumentation used during the observational part of the thesis, from the satellites and their instruments to the telescopes and their filters.

Chapters 4 and 5 contain the observational study of the two different flares studied in Ly$\alpha$ which were observed on 08 September 1999 and 28 February 1999 and the one observed in H$\alpha$ on 03 July 2002.
The description of the radiative transfer code used during the thesis and the discussion of the results is reported in Chapter 6.

Chapter 7 details the results of the Radiative MHD code applied to solar flare and the comparison of the Ly$\alpha$ and H$\alpha$ intensities with the observations and the radiative transfer results.

And finally Chapter 8 summarises the main results in this thesis and provides some ideas for a future work.

1.2 The layers of the Sun

The Sun is a G2V star located along the main sequence, with 70% of its mass composed of hydrogen, 28% helium and the rest made up of heavier elements. Its luminosity has been maintained during the last 4.6 billion years and it is calculated to live for another five billion years. Its mass is $1.99 \times 10^{30}$ kg and its effective temperature at the surface is about 5778 K. It has a mean diameter of $1.39 \times 10^9$ km, a luminosity of about $3.85 \times 10^{26}$ W and an apparent magnitude of -26.74. (Stix 2004). Table 1.1 shows very briefly the most important properties of the Sun.

The layers of the Sun are (See Fig. 1.1(a) for a schematic example):

1. The core: It extends about 0.25 $R_\odot$. The energy is created here by nuclear fusion, mainly of hydrogen by the p-p chain. Here the temperature is about $15.7 \times 10^6$ K and the density is up to 150 g/cm$^3$.

2. The radiative zone: It goes from 0.25 to 0.7 $R_\odot$. The energy generated in the core is transferred by electromagnetic radiation. A photon takes an average of 171 thousand years to leave the radiative zone.

3. The convective zone: It ranges from 0.7 to 1.0 $R_\odot$. Here the energy is transferred through the process of convection, more efficiently than in the radiative zone.

4. The atmosphere. It is divided into photosphere (the innermost layer), chromosphere, transition region, corona and heliosphere, which ends with the heliopause, where the interestellar medium and solar wind pressures balance (See Fig. 1.1(b)).

1.2.1 The Solar Atmosphere

The solar atmosphere is deemed to be part of the Sun layers above the visible surface, the photosphere, the surface that we can see with our naked eye or with optical telescopes. At this layer the temperature is about 5800 K.

<table>
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<td>Temperature Surface</td>
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<tr>
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Table 1.1: Some fundamental physical characteristics for the Sun.
1.2. THE LAYERS OF THE SUN

Above the photosphere, the temperature starts to rise even as the density continues to drop. The chromosphere, a layer of few hundred kilometers thick is as hot as 20000 K whereas the corona, even further up, is 1-5 MK. Figure 1.2 illustrates how the temperature and density profiles change as a function of height above the photosphere and the locations at which various emission lines are formed.

1.2.1.1 The Photosphere

The photosphere is defined as the inner surface of the Sun as seen in white light. It is only about 500 km thick and represents the boundary at which plasma becomes optically thick (i.e. the optical depth at 5000 Å = 1). The source of this opacity is due to the absorption of visible light by abundant negatively charged hydrogen ($H^-$) in the process of photoionisation (Wildt 1938). The surplus electron is easily removed by the hydrogen atom, in a bound-free continuum emission, emitting low-energy photons. The abrupt change in opacity is due to the decrease in $H^-$.

A number of features can be observed in the photosphere: the dark sunspots, the bright faculae and granules.

The most striking feature of the photosphere is the presence of sunspots (Fig. 1.3(a)): these dark regions consist of two parts: the umbra, which is the dark core of the spot, and the penumbra (“almost shadow”), which surrounds it. These structures, cooler than the surrounding material ($\approx 4000$ K compared to $\approx 6000$ K), are due to the presence of intense magnetic fields ($\approx 2500$-$3000$ G). Sunspots, together with the plages, can be considered the base of overlying active regions, the field lines of which are rooted in the photosphere.

The granules (Fig. 1.3(b)) have a size of the order of 1000 km, lifetimes of the order of 5 minutes and a mean vertical velocity lower than 0.5 km/s (Zirin 1966).

Solar faculae (Fig. 1.3(c)) are bright areas located near sunspots or in polar regions. They have sizes of 0.25 arcsec and a life duration between 5 minutes and 5 days. They are well seen in white light near the solar limb and with the help of

Figure 1.1: Schematic representation of the Sun. (a): The layers of the Sun; (b): Various features in the solar atmosphere. Pictures taken from http://www.windows2universe.org/sun/Solar_interior/Sun_layers/Convection_zone/sun_radzone_big_jpg_image.html and http://solar.physics.montana.edu/YPOP/Spotlight/Tour/images/sun_label.jpg
1.2.1.2 The Chromosphere

The chromosphere is the narrow layer (≈ 2500 km) of the solar atmosphere just above the temperature minimum of the photosphere and is usually visible as a thin red ring during a total solar eclipse. It is from this feature that it originally obtained its name: literally “colour-sphere” (its redness is due to the strong $H\alpha$ line). (Bray & Loughhead [1974].)

In the chromosphere we can distinguish:

- The Chromospheric Network: is a pattern seen in the red line of hydrogen ($H\alpha$) and in the ultraviolet line of calcium (Ca II K) (See Fig. 1.4 (a)).

Figure 1.2: Temperature and density of the solar atmosphere as a function of height. The heights of formation of several lines are reported. Taken from Vernazza et al. (1981).

Figure 1.3: Some features observed in the photosphere. (a): Sunspot (Courtesy: Royal Swedish Academy of Sciences); (b): Solar granulation (Courtesy: T. Rimmlele); (c): Solar faculae (Image taken from http://www.daviddarling.info/images/facula.jpg).
1.2. THE LAYERS OF THE SUN

- **Plage**: They are bright patches around sunspots, seen in Hα (See Fig. 1.4(b)).

- **Spicules**: are small, high-speed jets-like eruptions seen all over the chromospheric network. They appear as short dark streaks in Hα (Fig. 1.4(c)). They last a few minutes, ejecting material for 5-10 minutes out of the surface and into the hot corona at speeds of 20-30 km/s.

- **Filaments**: are dark, thread-like features seen in Hα. They are dense and cooler material than the surrounding, suspended above the solar surface by magnetic arcades (See Fig. 1.4(b)).

- **Prominences**: Prominences and filaments are physically the same phenomenon, except that prominences are seen projecting out above the limb. Both filaments and prominences can remain in a quiescent state for days or weeks. However, as the magnetic loops that support them slowly change, filaments and prominences can erupt and rise off in a timescale between few minutes to hours (See Fig. 1.4(d)).

Dominant chromospheric Fraunhofer lines are the CaII, H and K lines. The CaII lines are controlled by collisions and are suited to determine the chromospheric temperature.

An understanding of the heating of the chromosphere moves us into the world of magnetohydrodynamics and the chromosphere provides a valuable testing-ground for many aspects, particularly in cases where radiation losses must be taken into account.

---

1 It is a French word that means beach.
1.2.1.3 The Transition Region

In Figure 1.2 we can see that the photosphere, chromosphere and corona have quite smooth temperature and density profiles; the transition region, however, appears almost as a discontinuity sandwiched between the chromosphere and the corona.

The temperature jumps from \( \approx 25,000 \text{ K} \) to over 1,000,000 K in less than 100 km, indicating the presence of an efficient non-thermal source of heating. The composition of the transition region also changes over this short distance, from predominantly hydrogen and helium just above the chromosphere, to highly ionized, less abundant heavier ions nearer the corona (see Stix (2004)).

1.2.1.4 The Corona

The corona is the outermost layer of the solar atmosphere, which extends out various solar radius, becoming the solar wind. In the optical, after \( \sim 3 R_\odot \), the received intensity is almost zero and it is used to say that it extends to \( 3 R_\odot \). The corona is seen as an extended tenuous halo during a total solar eclipse; until the advent of coronographs, eclipses were the only opportunity to observe the corona: in the visible range, is six orders of magnitude fainter than the photosphere.

During one eclipse in 1869, the discovery of an emission line at 5303 Å incited to propose a new element, the “coronium”. More than 70 years later, this line was identified as a transition of \( \text{Fe}^{XIV} \) ion, formed at \( 2 \times 10^6 \text{ K} \) (Edlén 1943). The fact that the corona is a hundred or even a thousand times hotter than the photosphere poses the corona heating problem. When considering this problem one must identify an energy source which, when released, can heat the surrounding plasma taking also into account the response of the solar atmosphere. There are two different branches to explain the energy storage:

1. Energy stored in stressed magnetic field (DC heating).

2. Energy stored in plasma wave oscillations (AC heating).

A full discussion of the coronal heating problem, DC and AC heating mechanisms is discussed in a review by Klimchuk (2006) and references therein.

Coronal emission consists of two separate continuum spectra with a superimposed emission-line spectrum. (See Stix (2004) and Bray & Loughhead (1974)).

- The K-corona (\textit{Kontinuierlich}) is found at heights \( h \leq 0.3R_\odot \) and it is due to the electron scattering of photospheric continuum light.

- The L-corona, which arrives at heights \( h \leq 0.5R_\odot \), is dominated by line emission from highly ionised ions.

- The F-corona (\textit{Fraunhofer}) \( (h \geq 0.5R_\odot) \) is due to photospheric light scattered by dust particles, but also features the Fraunhofer absorption lines.

There are two types of coronal structures: those with open magnetic field lines and those with closed magnetic field lines.

1. Open-field regions, known as coronal holes, essentially exist at the solar poles (see Fig. 1.5(a)) and are the source of the fast solar wind \( (\approx 800 \text{ km s}^{-1}) \), which essentially moves plasma from the corona out into interplanetary space, appear darker in EUV and X-ray. (Parker 1958).
1.2. THE LAYERS OF THE SUN

Figure 1.5: Some features observed in the corona. (a): X-ray (from 6 to 49 Å) solar images from Skylab’s soft X-ray telescope; coronal holes are the dark regions where the hot coronal material is very thin. They were observed to rotate fairly rigidly and maintain their shape through several 27-day solar rotations; (b): Solar Corona in a false-color, 3-layer composite from the TRACE satellite: the blue ($1 \times 10^6$ degrees), green ($1.5 \times 10^6$ degrees), and red ($2 \times 10^6$ degrees) channels show the 171 Å, 195 Å and 284 Å wavelengths respectively. The corona has some (red) hot active regions in both hemispheres, surrounded by the (blue/green) cooler plasma of the quiet-Sun corona.

2. Closed magnetic field lines commonly form active regions (See Fig. 1.5(b)), which are the source of most of the explosive phenomena associated with the Sun.

At the minimum of the 11 years solar cycle, the corona is symmetric, with long streamers extending far out along the equator. At maximum, the corona is less symmetric about the equator, because there are many sunspots over the Sun, and from each active region there extends a long streamer.
Chapter 2

Solar Flares

Solar flares were defined by Smith & Smith (1963) as sudden, short lived brightening of a localized area of the chromosphere. They release vast quantities of energy \(10^{29} - 10^{32} \text{ ergs}\) on relatively short timescales (from a few minutes to tens of minutes). Nowadays we have a different view of flares: they are visible not only in the chromosphere, but also in the corona and photosphere.

Since the first confirmed observation of a solar flare in 1859 (Carrington 1859; Hodgson 1859), scientists have been fascinated about how this rapid release of energy occurs. They have been observed near large active regions with intense magnetic field and complex magnetic configuration, usually with an abrupt magnetic field gradient across the neutral line\(^1\), where the plasma is heated to temperatures higher than 15 MK (Régnier & Canfield 2006).

The energy release occurs at about 15,000 km, above the photosphere (Schrijver et al. 2008), emitting across the entire electromagnetic spectrum, from decametric radio emission to high-energy \(\gamma\) rays, heating the plasma to temperatures up to 20 MK. A sizeable fraction of this energy goes into accelerating particles (mainly electrons and protons) and nowadays it is not fully understood how the energy is released and converted into thermal heating and into nonthermal particle acceleration energy.

2.1 Flare Classification

Flares are ranked in importance depending on the intensity of their emission in optical, radio or X-ray. Table 2.1 shows the three types of flare classification from optical H\(\alpha\) images, radio data at 5 GHz and soft X-ray at 1-8 Å.

The most used is the soft X-ray one, according to the peak flux, measured by GOES (Geostationary Orbiting Environmental Satellite) at 1-8 Å: flares are classified into X, M, C, B and A; where each class has a peak flux ten times greater than the preceding one and X class flares have a peak flux of \(10^{-4} \text{ W m}^{-2} \equiv 10^{-1} \text{ erg cm}^{-2} \text{ s}^{-1}\). Within a class there is a linear scale from 1 to 9, so a C2 flare is twice as powerful as a C1 flare (e.g., \(M3 = 3 \times 10^{-2} \text{ erg cm}^{-2} \text{ s}^{-1}\)).

The H\(\alpha\) classification takes into account the area covered by a flare in the H\(\alpha\) images, from S, the smallest flares, to 4, the biggest ones (see Table 2.1). To give some information about the brightness of the flare, there is added a letter indication whether the intensity of the flare area is faint (f), normal (n), or brilliant (b): a Sb flare will be a brilliant flare with an area < 3 \(\times\) 10\(^8\) km.\(^1\)

\(^1\)A neutral line is defined as the line that separates solar magnetic fields of opposite polarity, typically determined from solar magnetograms recording the longitudinal magnetic component.
## 2.2 Typical Scenario of a Solar Flare

Many models have been elaborated to explain how a solar flare occurs. The current view based on the model of Petschek (1964) suggests that the magnetic energy is turned into plasma heating and accelerated electrons in the corona to energies of up to (10-100) keV. They follow the connected field lines to lower layers till the loop footpoints, usually situated at the chromosphere, where density is much higher (Kane 1983). The electrons that spiral along magnetic field lines produce microwave bursts via synchrotron radiation, losing their energy through Coulomb collisions and emitting hard X-ray emission via bremsstrahlung. This process is described by the thick target model (Brown 1971). See Section 2.7.2. This results in the so-called footpoint emission observed at hard X-ray energies (Hoyng et al. 1981; Sakao 1994; Petrosian et al. 2002). Accelerated ions, while colliding with background particles, can excite nuclear reactions and produce γ-ray emission (e.g. Ramaty & Murphy (1987); Hurford et al. (2003)).

In addition to the hard X-ray emission at the flare footpoints, observations often show also a third source located in the corona, at the top of the coronal loops, which presents a nonthermal component, associated with thin target emission of electrons propagating in a collisionless plasma (Datlowe & Lin 1973; Mariska & McTiernan 1999; Krucker & Lin 2008). In the thin target scenario, the electrons only lose a small fraction of their energy and continue towards the footpoints where they produce hard X-ray. Thin target emission is discussed in more detail in Section 2.7.1.

The released energy transferred to the lower atmosphere heat the chromosphere rapidly. The resulting overpressure in the overheated chromosphere can displace a mass flow upward along the loop at a speed of a few hundred km s−1. This mass motion fills the flaring loop with a hot plasma, giving rise to the loop structure seen in soft X-ray and gradual evolution of soft X-ray flux. This process was termed chromospheric evaporation by Neupert (1968) (See Chapter 2.6).

Energy redistribution in the lower layers of the atmosphere can produce ribbons seen in Hα, UV, EUV and occasionally in white-light (e.g. Hudson et al. (2006)) for the brightest flares. As time proceeds, reconnection develops to higher altitudes and the two hard X-ray footpoints and Hα ribbons are usually seen move away from each other, in a direction more or less parallel to the magnetic neutral line (Bogachev et al. 2005). This gives the standard picture of a two-ribbon flare (See Section 2.4.1).
2.3. SOLAR FLARE EVOLUTION

2.3 Solar Flare Evolution

The majority of flares can generally be described as having three phases (Kane 1974): one initial phase before the instability in the magnetic configuration called preflare phase and two distinct phases that occur after the reconnection process and the initial release of energy in the corona: an impulsive phase (or rise phase) and a gradual phase (or decay phase) (Kane & Anderson 1970). Figures 2.1 and 2.2 illustrate the profiles of these phases as observed in various energy ranges.

2.3.1 Preflare Phase

In the preflare phase the coronal plasma in the flare region slowly heats up and is visible in soft X-ray and EUV. The preflare hot thermal fluxes increase gradually for several minutes but are not detectable at energies above 10-20 keV ($T = 3 - 4 \times 10^7$ K).

2.3.2 Impulsive Phase

The observational situation preceding the impulsive phase is not cleared. Although a “precursor” soft X-ray brightening often occurs, it appears normally to be in a distinctly different location than the flare itself (Farnik & Savy 1998). Although observers note the gradual rise of a filament, small-scale brightenings, magnetic flux cancellations etc, there is no clear and unambiguous herald of the impulsive flare energy release.

In this phase most of the energy is released and a large number of energetic particles is accelerated, lasting from few tens seconds to few minutes (See Fig. 2.3). In the ‘standard flare model’, the accelerated particles propagate along the magnetic field lines to the chromosphere. These electrons exchange energy (i.e. heat) with surrounding electrons (equal mass interactions) but scatter and produce bremsstrahlung from...
their interaction with photons producing hard X-ray (HXR; 20-200 keV) emission that increases with several spikes (e.g., Tandberg-Hanssen & Emslie (1988); Zirin (1998)). The chromosphere responds dynamically through the process of chromospheric evaporation: as the plasma is heated to tens of MK, the thermal pressure becomes greater than that of the surrounding plasma. This creates a pressure gradient which drives the heated plasma through the flare loop into the corona, where the loop is seen emitting at soft X-ray and EUV wavelengths. The coronal thermal soft X-ray emission usually reaches its maximum after the impulsive phase.

During the impulsive phase, the coronal density increases by factors up to $10^3$ (Hudson & Ohki 1972). Later on, as the corona cools (Moore et al. 1980), Hα arcades form. Simultaneously, chromospheric Hα emission is concentrated into ribbons which slowly move apart.

### 2.3.3 Gradual Phase

In the decay or gradual phase, the coronal plasma returns almost to its original state after the gradual phase, except in the high corona ($\geq 1.2R_\odot$), where plasma ejections continue to accelerate particles causing interplanetary particle events. The effects on the corona become apparent in the soft X-rays and extreme UV loop arcades which form there.

The gradual phase may last several hours, depending on the magnitude of the
2.4 MORPHOLOGY OF SOLAR FLARES AND RELEVANT PHENOMENA

The morphology of solar flares is exceedingly diverse and complex, and no model accounts for all details. However, flares can roughly be divided into two groups (Pallavicini et al. 1977):

1. Two-ribbon flares.
2. Small, compact flares.

Figure 2.3: Impulsive phase of the 3 November 2003 X3.9 flare. (a): GOES 0.5-4 and 1.8 Å fluxes. The insert plots the GOES evolution from 9:00 to 15:00 UT, showing the pre-flare and gradual phases; (b)-(e): RHESSI light curves obtained in four energy bands with 4 second integration: 12-25, 25-50, 50-100 and 100-300 keV. (Veronig et al. 2006).

In Fig. 2.4(a), we see the light curve of an X1.5 flare which occurred on 6 March 1986 in different energy bands, where we can recognize:

- The pref flare phase, from about 13:50 to 13:56 UT, in which the soft X-ray emission gradually increases.

- The impulsive phase, from ~ 13:56 to 14:06 UT, where the hard X-ray and γ-ray emission rises rapidly with many spikes. The soft X-ray flux rises faster than in the previous phase.

- The gradual phase, that starts at about 14:06 UT, where the hard X-ray and γ-ray fluxes start to decay away more or less exponentially with a time constant of minutes. The soft X-ray flux continues to rise to a later peak and then it falls exponentially with a significant longer time constant, sometimes as long as several hours.

- After 14:10 UT a secondary phase of hard X-ray and γ-ray emission happens, in which the fluxes vary more gradually than during the impulsive phase. The soft X-ray flux continues to fall smoothly.

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2.4.1 Two-ribbon Flares

Often flares happen right after the sudden disappearance of a filament: the post-flare loops are situated along the filament (situated in the same region that the main neutral line of the magnetic field), forming a right angle with the long axis of the filament. The emission comes also from the base of the loops, forming two ribbons which lie on both sides of the location of the filament.

During the gradual phase, the two ribbons move away, typically at 10 km/s. The space between them is then usually filled with a transverse filamentary pattern seen in Hα (the post-flare loops). It is quite common, however, that after some time (hours or days) a new filament is formed at the place of the previous one, and even that another “homologous” flare erupts (see Section 2.4.3).

During the development of two-ribbon flares, the post-flare loops, when seen as loop prominences above the solar limb, exhibit material falling down from the apex along their legs and into the chromosphere.

Sometimes it is difficult to identify the two (or more) ribbons of a large flare, and at other times the ribbons exhibit exotic forms. Fig. 2.5 shows the evolution of an M3.9 flare occurred on 18 November 2003 in TRACE 1600 Å, where we can see the two ribbon emission.

2.4.2 Compact Flares

Compact flares have this name because they are small, versus two ribbons events, even if the flare is not compacted in a small region (Svestka 1986). A compact flare probably takes place in a small loop in the lower corona or an arch filament system and usually it is possible to observe small changes in its structure; the emission is largely confined to the plasma in the loop where it eventually dies away.

Subflares are the smallest of the compact class flares; they are short-lived, being

![Figure 2.4: Time profile at different energy ranges of an X1.5 flare occurred on 6 March 1989 (Dunphy & Chupp 1991).](image)
2.4. MORPHOLOGY OF SOLAR FLARES AND RELEVANT PHENOMENA

Figure 2.5: Temporal evolution of the M3.9 flare occurred on 18 Nov. 2003 showing flare ribbons in TRACE 1600 Å images: a) impulsive phase; b) time of RHESSI maximum; c) time of GOES maximum; d) decay phase. FOV: 320" × 250". (Image from Miklenic et al. (2009)).

only slightly brighter than an active plage.

2.4.3 Homologous Flares

It was already noticed by Waldmeier (1938) that in many active regions, flares have a tendency to recur in nearly identical form seen in Hα. Ellison et al. (1960) proposed the term homologous flares to describe those cases where flares occur repetitively in the same region, with essentially the same position relative to the spot group and with a common pattern of development. Woodgate et al. (1984) demonstrated that the “homologous flares” have a similar shape in the Hα and UV images, even during the impulsive phase.

2.4.4 Flares Seen at the Solar Limb

Limb flares permit us to examine flares in cross-section projection, and provide significant information about their shapes as well as their dimensions.

2.4.5 Filament Eruptions

Although the violent ejection of visible material is not an essential part of the flare process, many flares exhibit such behavior. The most common type of flare ejection is the filament eruption. Prominences at the limb usually appear to grow upward from the chromosphere out into the corona. After reaching maximum extension, the material is usually seen to descend into the chromosphere at its point of origin, along the original trajectory.

In Fig. 2.6 we can see different filament eruptions associated with solar flares. Fig. 2.6(a) shows an erupting filament associated with an M-class flare observed on 29 April 1999; the chromospheric counterpart of the flare are two ribbons, seen with the dark filament in between. Fig. 2.6(b) shows a spectacular quiescent filament eruption observed on 12 September 2000, associated with an M-class flare. After 20 minutes from the eruption (middle image), we can see a dark filament in the northwestern
CHAPTER 2. SOLAR FLARES

Figure 2.6: BBSO Hα images taken from Tripathi et al. (2009). (a): Erupting filament on 29 April 1998. The arrow locates the surviving filament in between two ribbons. (b): Partially erupting filament on 12 Sept 2000, marked by arrows; (c): Filament before eruption (left panel), associated with a ribbon flare during eruption (middle panel) and surviving filament in between two ribbons (right panel).

side. This provides evidence that the filament breaks in the middle towards its north-western end during eruption. Fig. 2.6(c) shows an erupting filament observed on 13 May 2003, associated with an M-class flare. The right image shows that a two-ribbons flare brackets the surviving filament (See Tripathi et al. (2009)).

2.5 Flare Models

During solar flares, the release of the energy stored in the twisted magnetic field configuration proceeds through a process termed magnetic reconnection: the magnetic field topology is fundamentally changed, breaking and reconfiguring the field lines. The different flare models are mainly distinguished by the initial magnetic topologies.

There are still some poorly understood questions like: the geometry of the energy release region, the mechanism that causes the dissipation of the stored energy or the physical process that causes the energy to be released suddenly after many hours of energy buildup without significant dissipation of the energy stored. Sturrock (1980) provided a detailed review of the various reconnection scenarios that existed till the end of the 1970s and most of them are still valid, like for instance the model of Gold & Hoyle (1960), which assumes that there is a self attraction between the magnetic loops carrying parallel currents, causing an energy release at their mutual
2.5. FLARE MODELS

Figure 2.7: Version of the standard 2D X-type reconnection model that includes the slow and fast shocks in the outflow region, the upward-ejected plasmoid, and the locations of the soft X-ray bright flare loops.

The most widely-used model for flares is the 2D magnetic reconnection standard model that evolves from the concepts of Carmichael (1964a), Sturrock (1966), Hirayama (1974), Kopp & Pneuman (1976), called the CSHKP model according to the initials of their names, which explains the gradual phase of solar flares.

This model fits the observations in hard X-rays, soft X-rays, Hα, and radio wavelengths: it suggests that the soft X-ray loop prominence system grows in time, increasing also the Hα ribbons; it appeals a reconnection site near the apex of the loop system, injecting the energy into the underlying loops. The standard model does not specify what urges on the initial magnetic system to become unstable; it explains single loops and two ribbons arcade geometries, but does not fit for quadrupolar flare loop interactions or 3D nullpoint topologies (Aschwanden (2005), Chapter 10).

Tsuneta (1996, 1997) and Shibata (1995) added to the model (Fig. 2.7) the heated plasma in the reconnection outflow, which produces hot ridges (\(T \approx 15 - 20\) MK) along the separatrices, sandwiching the soft X-ray loops that occupy the newly reconnected field lines, filled with chromospheric evaporated plasma. The fast shocks in the reconnection outflows collide with the previously reconnected field lines and may produce hot thermal (as well as non-thermal) hard X-ray sources above the top of the flare loops.

The model of Tsuneta et al. (1997) uses the simultaneous imaging observations in soft and hard X-ray and suggests that the reconnection site is situated near the apex of the loop system, which injects energy (e.g., as accelerated particles) into the underlying loops. This model explains, in a quantitative way, structure of flares, even if it does not fit perfectly with the observations.

It was recognized that it is necessary to address the physics of magnetic reconnection, and the contribution of accelerated electrons in the context of the observations. Nowadays there are also many observations that require 3D models, multiple flare

---

2The impulsive phase of solar flares cannot be treated as a 2D model.
loops, as well as tripolar, quadrupolar or 3D nullpoint magnetic topologies.

2.6 Chromospheric Evaporation

The footpoints of flare loops are based in the chromosphere where the accelerated particles lose their energy producing the HXR emission and heating rapidly the chromospheric plasma to tens of millions of degrees, expanding and filling the overlying corona (see Figs. 2.8 and 2.9). This process, known as chromospheric evaporation, has been studied for almost 40 years.

The concept of chromospheric evaporation has gained much support over the years through observations of blueshifted emission of high-temperature plasma, often temporally correlated with impulsive HXR burst (e.g. 300-400 km s$^{-1}$ upflows) (Doschek et al. 1980; Antonucci & Dennis 1983).

During the impulsive phase of a flare, the atmosphere responds to the injection of a flux of accelerated electrons (Fisher et al. 1985b; Abbett & Hawley 1999). For electron fluxes $<10^{10}$ erg cm$^{-2}$ s$^{-1}$, the evaporated plasma flows upward at several tens of kilometers per second. This “gentle” evaporation due to non-thermal electrons has been observed by Milligan et al. (2006b).

At high non-thermal electron fluxes ($>3\times10^{10}$ erg cm$^{-2}$ s$^{-1}$), the chromosphere is unable to radiate at a sufficient rate and consequently it expands rapidly. This condition is met when the heating rate time-scale is less than the hydrodynamic expansion time-scale:

$$\frac{3kT}{Q} > \frac{L_0}{c_s}$$  \hspace{1cm} (2.1)

where $Q$ is the flare heating per particle, $T$ is the final temperature of the heated plasma, $c_s$ is the corresponding sound speed, and $L_0$ is the length-scale of the flaring region. Once this condition is achieved, the heated chromospheric plasma expands.
2.7. HARD X-RAY EMISSION AND FLARE SPECTRA

Figure 2.9: Chromospheric evaporation is caused by heating of precipitating electrons and nuclear $\gamma-$ rays \cite{Aschwanden2005}.

upward at hundreds of kilometers per second in a process known as “explosive” evaporation. The overpressure of the flare plasma relative to the underlying chromosphere causes cooler, more dense material to recoil downward at tens of kilometers per second. \cite{Milliganetal2006a} confirmed these flow patterns and that the upflowing plasma originates within the HXR footpoints.

2.7 Hard X-ray Emission and Flare Spectra

The X-ray flux spectrum of photons of energy $\epsilon$ due to electrons of energy $E$, assuming isotropic emission, is given by Eq. \ref{eq:2.2}:

$$I(\epsilon) = \frac{1}{4\pi R^2} \int_V \int_{\epsilon}^{\infty} n(\vec{r}) F(E, \vec{r}) Q(\epsilon, E) dEdV$$ \hspace{1cm} \text{(2.2)}

where $R$ is the distance to the X-ray detector; $Q(\epsilon, E)$, the bremsstrahlung cross section and $F(\vec{r}, E)$ (electrons $\text{cm}^{-2} \text{s}^{-1} \text{keV}^{-1}$), the electron beam flux density distribution.

In a dense plasma the bremsstrahlung losses are dominated by collisional losses to the plasma electrons. For a fully ionized plasma, the (nonrelativistic) loss rate is

$$\frac{dE}{dt} = -\frac{\kappa}{E} n(\vec{r}) \nu(E),$$ \hspace{1cm} \text{(2.3)}

where $\kappa = 2\pi e^4 \Lambda_{ee} \approx 2.6 \times 10^{-18} \text{ cm}^2 \text{ keV}^2$, $\Lambda_{ee}$ is the Coulomb logarithm, and $\nu(E)$ is the speed of the electron \cite{Brown1971}.

$$\nu dt = dz \rightarrow \frac{dE}{dN} = -\frac{\kappa}{E},$$ \hspace{1cm} \text{(2.4)}

where $N(z)$ (\text{cm}^{-2}) is the column density.

From Eq. \ref{eq:2.4} we get that the evolution of the energy of an electron with the column density is:

$$E^2 = E_0^2 - 2\kappa N,$$ \hspace{1cm} \text{(2.5)}

where $E_0$ is the initial (injected) energy of the electron.

If energy losses are not significant within an X-ray source, the emission is called thin-target; otherwise the emission is called thick-target.
2.7.1 Thin Target Model

For a thin target model the electrons are injected into a collisionless plasma where they undergo few collisions, leaving the injected distribution mostly unchanged.

The thin target photon flux measured at Earth can be related to the injected electron spectrum by (Tandberg-Hanssen & Emslie 1988):

\[ I_{\text{thin}}(\epsilon) = \frac{S \Delta N}{4\pi R^2} \int_{E_0}^{\infty} F(E_0) \sigma_B(\epsilon, E) dE_0 \]  

(2.6)

where \( R = 1 \) AU and \( \Delta N = \int_{\text{source}} n(s) ds \) is the column density of the source and \( n \) is the ambient target density. \( \Delta N \) will be less than the column depth required to completely stop an electron of energy \( E \).

2.7.2 Thick Target Model

In the thick target model electrons precipitate or thermalise in a collisional plasma due to Coulomb collisions.

Calculating the rate at which the electrons of energy \( E \) radiates bremsstrahlung photons of energy \( \epsilon \), the thick-target X-ray spectrum is then given by Eq. 2.7:

\[ I_{\text{thick}}(\epsilon) = \frac{1}{4\pi R^2} \int_{E_0=\epsilon}^{\infty} E Q(\epsilon, E) dE dE_0. \]  

(2.7)

Observed X-ray spectra from solar flares can usually be well fitted with a single power law electron flux distribution of the form \( F(E) \propto E^{-\delta} \). The photon spectrum also has the power-law form \( I(\epsilon) \propto \epsilon^{-\gamma} \). The relationship between the electron and photon spectral indices \( \delta \) and \( \gamma \) can be obtained using the Kramers approximation to the bremsstrahlung cross section: \( Q(\epsilon, E) \propto 1/\epsilon E \).

For a simple thin-target source,

\[ I_{\text{thin}}(\epsilon) \propto \epsilon^{-(\delta+1)}, \]  

(2.8)
giving \( \gamma_{\text{thin}} = \delta + 1 \).

For a thick-target source region,

\[ I_{\text{thick}}(\epsilon) \propto \epsilon^{-(\delta-1)}, \]  

(2.9)
giving \( \gamma_{\text{thick}} = \delta - 1 \).

Analytic expressions relating the normalization coefficients can also be obtained when the non-relativistic Bethe-Heitler bremsstrahlung cross section is valid (Brown 1971; Tandberg-Hanssen & Emslie 1988).

These simple power-law relationships are not valid if there is a break or a cutoff in the electron distribution above the energies of interest. For example, these relationships are not correct for the lower power-law index of a double power-law fit to a photon spectrum.

2.7.2.1 Energy cut-off in nonthermal electrons

For the thick-target model, the photon spectrum \( I(\epsilon) \) is directly determined by the injected electron flux distribution \( F_0(E_0) \). By integrating over all electron energies, we can also determine the total flux of nonthermal electrons, \( N_{\text{nth}} \) electrons \( (s^{-1}) \), the power in nonthermal electrons, \( P_{\text{nth}} \) (erg s\(^{-1}\)), and, integrating over time, the total energy in nonthermal electrons.
2.7. HARD X-RAY EMISSION AND FLARE SPECTRA

The non-thermal power is computed as follow:

\[ P_{nth} = \kappa_E \int_{E_c}^{+\infty} E_0 \mathcal{F}_0(E_0) dE_0 = \kappa_E A \frac{E_c^{-\delta+2}}{\delta - 2}, \]  

(2.10)

where \( E_c^{-\delta+2} \) is the result for a power-law electron flux distribution of the form \( \mathcal{F}_0(E_0) = A E_0^{-\delta} \); \( E_c \) is a low-energy cutoff to the electron flux distribution and the constant \( \kappa_E = 1.6 \times 10^{-9} \) is the conversion from keV to erg.

The non-thermal power (and the non-thermal energy) from the power-law electron flux distribution depends on three parameters: \( \delta, A \) and \( E_c \). Observations indicate that \( \delta \) is almost always greater than 2 (Dennis 1985; Lin & Schwartz 1987; Winglee et al. 1991; Holman et al. 2003). \( E_c \) cannot be null because the integral would yield an infinite value, a decidedly unphysical result.

The essence of the problem in many flare spectra is summarized in Fig. 2.10: the nonthermal power-law is well observed above 20 keV, but any revealing features that it might possess at low energies, such as a low-energy cutoff, are destroyed by the presence of the thermal emission.

The hard X-ray spectrum observed during a typical solar is shown in Fig. 2.10(a), where the black line represents the observed photons and crosses the preflare background. The models fitted are shown by colored lines. Fig. 2.10(b) shows the model spectra only.

The components shown in Fig. 2.10(b) are:

- Thermal: the purple line is the thermal emission associated to the hot plasma, formed by the continuum and the atomic lines. The black curve is what RHESSI should be able to see: the same spectrum seen with a resolution of 1 keV.

Figure 2.10: (a): Hard X-ray spectrum of a solar flare. This spectrum is in count space, therefore the model components have been folded with the instrumental response function; (b): Typical full sun flare spectrum. Dashed: Nonthermal spectrum from an accelerated electron distribution with \( \delta = 4 \), and a low-energy cutoff of 20 keV. Dotted: Thermal spectrum, from a plasma with temperature \( T=20 \) MK and emission measure \( EM = 10^{49} \text{ cm}^{-3} \). Solid: Total radiated spectrum. Images taken from Grigis (2006).
• Nonthermal: the red line represents a broken power-law function:

\[
F(E) = \begin{cases} 
F_0 \left( \frac{E}{E_0} \right)^{-\gamma_1} & \text{if } E \leq E_B \\
F_0 \left( \frac{E}{E_0} \right)^{-\gamma_1} \left( \frac{E}{E_B} \right)^{-\gamma_2} & \text{if } E > E_B
\end{cases}
\]

where \( \gamma_1 \) and \( \gamma_2 \) are the spectral indices below and above the break, respectively.

## 2.8 Magnetic Field

Flares are intimately connected with solar magnetic fields and the magnetic energy released after the reconnection of the field lines.

To have an idea about the magnetic field at the corona, MDI onboard SOHO measures the line-of-sight photospheric magnetograms, which is used as a boundary condition for numerical calculation of the coronal magnetic field.

The magnetic field extrapolation methods are based in various assumptions (Neukirch 2005):

• The coronal magnetic field is in equilibrium without flows.

• The ratio of thermal pressure to magnetic pressure (the plasma \( \beta \)) in the corona is small.

• The coronal structures change on length scales comparable to the typical coronal scale height.

### 2.8.1 Basic equations

The magnetized plasma of the solar atmosphere is usually well described by the equations of ideal magnetohydrodynamics given by:

\[
\frac{\partial \rho}{\partial t} + \nabla \cdot (\rho \vec{v}) = 0 \quad (2.11)
\]

\[
\rho \left( \frac{\partial \vec{v}}{\partial t} + \vec{v} \cdot \nabla \vec{v} \right) + \nabla p - \vec{j} \times \vec{B} + \rho \vec{g} = 0 \quad (2.12)
\]

\[
\nabla \times \vec{B} - \mu_0 \vec{j} = 0 \quad (2.13)
\]

\[
\frac{\partial \vec{B}}{\partial t} + \nabla \times \vec{E} = 0 \quad (2.14)
\]

\[
\vec{E} + \vec{v} \times \vec{B} = 0 \quad (2.15)
\]

\[
\nabla \cdot \vec{B} = 0 \quad (2.16)
\]

where \( \vec{B} \) is the magnetic field; \( \rho \), the plasma density; \( \vec{v} \), the plasma velocity; \( p \), the plasma pressure; \( \vec{j} \) is the current density; \( \vec{g} \), the gravitational acceleration; \( \mu_0 \), the vacuum permeability and \( \vec{E} \) is the electric field.

Assuming that the plasma is in equilibrium \( \frac{\partial \vec{v}}{\partial t} = 0 \) and that the effect of plasma flows can be neglected (\( \vec{v} \) small), the previous equations can be written in a more simplified way (Neukirch 2005), obtaining the equations for force-free equilibria:

\[
\vec{j} \times \vec{B} = \vec{0} \quad (2.17)
\]
2.8. MAGNETIC FIELD

\[ \nabla \times \vec{B} = \mu_0 \vec{j} \quad \text{(2.18)} \]
\[ \nabla \cdot \vec{B} = 0 \quad \text{(2.19)} \]

Taking into account that \( \vec{j} \times \vec{B} = 0 \), we can get from these equations that:
\[ \nabla \times \vec{B} = \alpha(\vec{r}) \vec{B} \quad \text{(2.20)} \]

and using the Eqs. (2.18), (2.19) and (2.20) we arrive to:
\[ \vec{B} \cdot \nabla \alpha = 0 \quad \text{(2.21)} \]

implying that the function \( \alpha \) is constant along magnetic field lines.

According to the value of \( \alpha \), we can distinguish three different force-free fields:

1. The simplest assumption is \( \alpha = 0 \) everywhere. By Eq. (2.20) this corresponds to potential fields.

2. If \( \alpha \) is assumed to be constant, but non-zero, we get the linear force-free fields.

3. The general case is when \( \alpha \) is not constant from field line to field line. It is called nonlinear force free case.
Chapter 3

Instrumentation

3.1 TRACE

The TRACE telescope (Transition Region and Coronal Explorer) was a NASA space telescope designed to get high resolution images of the solar photosphere and transition region to the corona. It was a small explorer mission, launched in April 1998 on a Pegasus rocket. On 21 June 2010 it stopped to work.

TRACE is being used to study the solar atmosphere: the corona, transition region and chromosphere conditions, the structure of coronal loops, the evolution of the different features observed, etc.

The primary and secondary mirrors of the TRACE telescope are divided into quadrants, and a rotating shutter selects the quadrant that will be illuminated (see Fig. 3.1) (Handy et al. 1999a). Three of the quadrants are coated with multilayers for imaging at EUV wavelengths, and the fourth quadrant is coated with aluminium and magnesium fluoride for imaging very broad ultraviolet wavelength ranges near 1216, 1550, 1600 and 1700 Å. The different wavelength passbands correspond to plasma emission temperatures from 4,000 to 4,000,000 K. (See Table 3.1). Images in all the wavelengths are projected onto a single 1024 x 1024 CCD detector with a spatial resolution of 0.5 arcsec/pixel on a side, giving an 8.5 arcmin field of view.

Figure 3.1: The TRACE telescope.
### Table 3.1: TRACE temperature response

<table>
<thead>
<tr>
<th>Wavelength (Å)</th>
<th>Emission</th>
<th>Bandwidth (Å)</th>
<th>Temperature (K)</th>
</tr>
</thead>
<tbody>
<tr>
<td>171</td>
<td>Fe IX/X</td>
<td>6.4</td>
<td>1.6 - 2.0 \times 10^5</td>
</tr>
<tr>
<td>195</td>
<td>Fe XII/XXIV</td>
<td>6.5</td>
<td>5.0 - 20 \times 10^5</td>
</tr>
<tr>
<td>284</td>
<td>Fe XV</td>
<td>10.7</td>
<td>1.25 - 4.0 \times 10^6</td>
</tr>
<tr>
<td>1216</td>
<td>HI Lyα</td>
<td>84</td>
<td>1.0 - 3.0 \times 10^4</td>
</tr>
<tr>
<td>1550</td>
<td>C IV</td>
<td>30</td>
<td>6.0 - 25 \times 10^4</td>
</tr>
<tr>
<td>1600</td>
<td>UV Cont, C I, Fe II</td>
<td>275</td>
<td>4.0 - 10 \times 10^3</td>
</tr>
<tr>
<td>1700</td>
<td>Continuum</td>
<td>200</td>
<td>4.0 - 10 \times 10^3</td>
</tr>
<tr>
<td>5000</td>
<td>White Light</td>
<td>Broad</td>
<td>4.0 - 10 \times 10^3</td>
</tr>
</tbody>
</table>

3.2 SOHO

The SOlar and Heliospheric Observatory (SOHO), is a project of ESA and NASA. It was launched on December 2, 1995 and with the 12 instruments it was onboard, it has studied the Sun from its interior, through the atmosphere, to the solar wind.

Some of the key results include ([http://sci.esa.int/science-e/www/object/index.cfm?fobjectid=38381](http://sci.esa.int/science-e/www/object/index.cfm?fobjectid=38381)):

- Studying the convection zone and the sunspots below the surface.
- Providing information about the temperature structure, the interior rotation, and plasma flows in the solar interior.
- Measuring the slow and fast solar wind.
- Finding the source of the fast solar wind in the magnetically open regions.
- Forecasting the space weather, giving three days notice of the disturbances at the Earth.

The satellite has two main elements: the payload module, boarding the 12 instrument packages, and the service module, providing essentials such as power and communications (See Fig. 3.2).

During my Ph.D. I used mainly data from the Michelson Doppler Imager (MDI) and Extreme ultraviolet Imaging Telescope (EIT) instruments, which are detailed in Sections 3.2.1 and 3.2.2.

3.2.1 MDI

The Michelson Doppler Imager (MDI) is a part of the Solar Oscillations Investigation (SOI) instrument, on board SOHO (See Fig. 3.3(a)).

It images the Sun on a CCD camera (1024 x 1024 pixels). MDI is able to measure dopplergrams, line-of-sight magnetograms, intensity, line-depth, and four filtergrams or white light intensity.

A pair of Michelson interferometers, enable MDI to record filtergrams with a FWHM bandwidth of 100 mÅ. The filtergrams are centered at four wavelengths equally spaced by 75 mÅ across the NiI 6768 Å spectral line; they are obtained each minute and used to calculate the velocity and continuum intensity with a resolution of 4 arcsec over the whole disk ([Scherrer et al., 1995](http://www.solarviews.com)).
A line-of-sight magnetogram is determined using the difference of two Doppler-grams from left and right circular polarization. There are full disk and high resolution magnetograms:

- Full disk images: with a field of view of $34 \times 34$ arcmin$^2$ with a resolution of 1.98 arcsec/pixel.

- High resolution images: with a field of view of $11 \times 11$ arcmin$^2$ with spatial resolution of 0.62 arcsec/pixel.

Normally MDI produces full disk magnetograms in a cadence of 96 minutes.

### 3.2.2 EIT

The Extreme ultraviolet Imaging Telescope (EIT) is one of the 12 instruments on board SOHO (See Fig. 3.3 (b)). It provides full disk images of the corona and transition region on the solar atmosphere. It observes at four selected wavelengths in the extreme ultraviolet: Fe IX (171 Å), Fe XII (195 Å), Fe XV (284 Å) and He II
(304 Å), providing a temperature diagnostics in the range from $6 \times 10^4$ K to $3 \times 10^6$ K (See Table 3.2).

The telescope has a $45 \times 45$ arcmin$^2$ field of view and spatial resolution of 2.6 arcsec/pixels.

### 3.3 Yohkoh

The Yohkoh satellite (Fig. 3.4) is a project of the Institute for Space and Astronautical Sciences. It studied for ten years X-rays and $\gamma$-rays emission from the Sun. Yohkoh was launched from Kagoshima, Japan on 31 August 1991, it was lost on 14 December 2001 and on 12 September 2005 the spacecraft burned up during reentry over South Asia. The name Yohkoh is Japanese for “sunbeam”.

Onboard the satellite there were four instruments that detected energetic emissions from the Sun (http://umbra.nascom.nasa.gov/yohkoh/docs/yag/iguide/iguide.html):

- the Bragg Crystal Spectrometer (BCS)
- the Wide Band Spectrometer (WBS)
- the Soft X-Ray Telescope (SXT)
- the Hard X-Ray Telescope (HXT).

The Yohkoh Spacecraft was in a slightly elliptical low-earth orbit, with an altitude ranging from approximately 570 km to 730 km. The orbital period was 90 minutes (65-75 minutes of this time was spent in sunlight).

<table>
<thead>
<tr>
<th>Wavelength</th>
<th>Ion</th>
<th>Peak Temperature</th>
<th>Observational Objective</th>
</tr>
</thead>
<tbody>
<tr>
<td>304 Å</td>
<td>He II</td>
<td>$8.0 \times 10^4$ K</td>
<td>chromospheric network; coronal holes</td>
</tr>
<tr>
<td>171 Å</td>
<td>Fe IX-X</td>
<td>$1.3 \times 10^6$ K</td>
<td>corona/transition region boundary; structures inside coronal holes</td>
</tr>
<tr>
<td>195 Å</td>
<td>Fe XII</td>
<td>$1.6 \times 10^6$ K</td>
<td>quiet corona</td>
</tr>
<tr>
<td>284 Å</td>
<td>Fe XV</td>
<td>$2.0 \times 10^6$ K</td>
<td>active regions</td>
</tr>
</tbody>
</table>

Table 3.2: EIT Bandpasses Delaboudinière et al. (1995).
3.3. YOHKOH

3.3.1 SXT

<table>
<thead>
<tr>
<th>Filter</th>
<th>Centered Wavelength</th>
<th>Wavelength Range</th>
</tr>
</thead>
<tbody>
<tr>
<td>Open</td>
<td>no filter analysis</td>
<td>2.5 - 46 Å</td>
</tr>
<tr>
<td>Al 1265 Å</td>
<td>1265 Å</td>
<td>2.5 - 36 Å</td>
</tr>
<tr>
<td>Dagwood Sandwich</td>
<td>2930 Å Al 2</td>
<td>2.4 - 32 Å</td>
</tr>
<tr>
<td></td>
<td>2070 Å Mg 562 Å Mn</td>
<td></td>
</tr>
<tr>
<td></td>
<td>190 Å C</td>
<td></td>
</tr>
<tr>
<td>Be 119 mm</td>
<td>119 µm</td>
<td>2.3 - 10 Å</td>
</tr>
<tr>
<td>Al 12 mm</td>
<td>11.6 µm</td>
<td>2.4 - 13 Å</td>
</tr>
<tr>
<td>Mg 3 mm</td>
<td>12.52 µm</td>
<td>2.4 - 23 Å</td>
</tr>
<tr>
<td>Wide band optical filter</td>
<td></td>
<td>4600 - 4800 Å</td>
</tr>
<tr>
<td>Narrow band optical filter</td>
<td></td>
<td>4290 - 4320 Å</td>
</tr>
</tbody>
</table>

Table 3.3: Yohkoh/SXT filters (http://umbra.nascom.nasa.gov/yohkoh/docs/ydac/rguide/subsection3_3_4.html).

The soft X-ray telescope (SXT) on Yohkoh [Tsuneta et al. 1991] has a spatial resolution of 2.5 arcsec/pixel. The flare images have a temporal resolution of 2 seconds, while the typical time resolution in the quiet mode is 8.0 seconds.

The SXT imaged X-rays in the 0.25 - 4.0 keV range. It used thin metallic filters to acquire images in restricted portions of this energy range.

The characteristics of the SXT filters are reported in Table 3.3. After the failure of an entrance filter in November 1992, the narrow band, wide band optical filters, and Noback X-ray filter became unusable.

3.3.2 HXT

The Yohkoh hard X-ray telescope (HXT, [Kosugi et al. 1991]) uses a Fourier-synthesis technique to take images in four hard X-ray (HXR) energy bands (http://umbra.nascom.nasa.gov/yohkoh/docs/ydac/rguide/subsection3_3_3.html); i.e., L (14-23 keV), M1 (23-33 keV), M2 (33-53 keV) and H (53-93 keV) through sixty-four pairs of grids (See Table 3.4), being able to reconstruct an image of the source in each of the four energy bands.

The field of view of HXT is about 35 × 35 arcsec², covering the whole Sun. The spatial resolution is 5 arcsec/pixel. The temporal resolution is up to 0.5 seconds [Kosugi et al. 1991].

<table>
<thead>
<tr>
<th>Energy Band</th>
<th>Energy Range</th>
<th>Background count rate</th>
</tr>
</thead>
<tbody>
<tr>
<td>L</td>
<td>13.9 - 22.7 keV</td>
<td>≈ 1 cts/s/SC</td>
</tr>
<tr>
<td>M1</td>
<td>22.7 - 32.7 keV</td>
<td>≈ 2 cts/s/SC</td>
</tr>
<tr>
<td>M2</td>
<td>32.7 - 52.7 keV</td>
<td>≈ 1 cts/s/SC</td>
</tr>
<tr>
<td>H</td>
<td>52.7 - 92.8 keV</td>
<td>≈ 8 cts/s/SC</td>
</tr>
</tbody>
</table>

Table 3.4: Yohkoh/HXT energy bands.
3.4 RHESSI

The (Reuven) Ramaty High Energy Solar Spectroscopic Imager (RHESSI, Fig. 3.5) is one of NASA Small Explorer missions (Lin et al. 2002), launched in February 2002. It studies mainly the particle acceleration and explosive energy release in solar flares.

The single instrument is designed to observe energies from $3$ keV (soft X-ray) to $17$ MeV ($\gamma$-rays), with an energy resolution of about $1$ keV FWHM in the hard X-ray range and a few keV in the $\gamma$-ray range. RHESSI has 9 identical germanium detectors with a diameter of 7.1 cm each and performs imaging and spectral analysis; in front of each detector there is a pair of grids: the rotation of the satellite (about 15 RPM) produces a modulation of the signal of sources seen through the grids with a spatial resolution of $2.3$ arcsec/pixel ($\approx 1660$ km).

For large count rates, two shutters can be put in place giving three attenuator states, A0 (open state), A1 (thin attenuator) and A3 (thick attenuator). The effect of the attenuator state is energy dependent and is more significant at lower energies where the counts are highest, allowing higher energy photons to penetrate.

Figure 3.5: The Ramaty High Energy Solar Spectroscopic Imager (RHESSI) spacecraft.

Figure 3.6: (a): Schematic view showing RHESSI Rotating Modulation Collimators Hurford et al. (2002); (b): Schematic representation showing the imaging process.
3.4. RHESSI

3.4.1 Imaging

The instrument has 9 rotating modulating collimators; each collimator consists of two grids separated by 1.5 meters and made by an array of opaque and transparent equally spaced slits (Fig. 3.6). The spatial resolution depends on the slit width, reaching angular resolution down to 2.3 arcsec.

Behind each collimator there is the detector, which records the arrival, time and energy of the photons (Smith et al. 2002), which is modulated due to the rotation of the spacecraft.

A combination of the time-modulated flux from different detectors provides information required for effective image construction. Fig. 3.7 shows examples of the time-modulated flux for an off-axis source with varying parameters:

- Panel 1 shows the modulation pattern for one rotation for an off-axis point source.

- Panel 2: the same point source but half the flux thus reducing the amplitude of the modulated signal by half.

- Panel 3: another point source, at the same radial distance from the rotation axis centre, but at a different position on the Sun. The effect of this is a phase shift in the modulated signal.

- Panel 4: a point source at a greater radial distance from the rotation centre. This is reflected in the frequency of the modulated pattern as the slit passes across the source in a shorter time.

Figure 3.7: Examples of modulation profiles for an off-axis source. Different panels correspond to different source parameters such as radial offset and source size. Figure taken from (Lin et al. 2003).
CHAPTER 3. INSTRUMENTATION

- Panel 5: a source with finite size which results in the trough of each period occurring at a finite value instead of the signal dropping to zero.
- Panel 6: an even larger source size which enhances the effect shown in panel 5.
- Panel 7: a more realistic modulation pattern which contains a mixture of all the previously mentioned features.

It is possible to reconstruct an image from the time-modulated signal using an inverse Fourier transform, but this process is affected by the instrumental response. A number of algorithms have been developed to remove these instrumental effects and return the true source. A full discussion of these algorithms is given in Smith et al. (2002).

3.4.2 Spectroscopy

For the spectral analysis of the RHESSI data, there is a program (OSPEX, (Schwartz et al. 2002)) which fits the counts measured by the detector to a spectrum model.

![Figure 3.8: A cutaway of the RHESSI Spectrometer showing the location of the detectors under each grid (by number) (Lin et al. 2002).](image)

3.5 Ondřejov Solar Observatory

The Solar Ondřejov Telescope is a horizontal one with the main objective mirror of 230 mm/1350 cm. Fig. 3.9 shows the telescope and spectrograph systems. The horizontal solar telescope creates the image on the slit of the spectrograph (See Fig. 3.9).

In the case of photographic observations, usually the exposure time is about 0.1 - 0.3 seconds for flares situated at the solar disk and 10 seconds for prominences, when the second diffractional order is used.

3.5.1 Multichannel Flare Spectrograph

The Multichannel Flare Spectrograph (MFS) was built in 1958 (Valněček et al. 1959) and was used for detection of photographic spectra in several diagnostically important lines simultaneously for more than 2 solar cycles, until June 2004.

The spectra are used mainly for quantitative plasma diagnostics of the flaring atmosphere or prominence structures, as well as of the dynamic processes like velocity fields. After several reconstructions (Kotrč et al. 1993, Kotrč 1997), the instrument...
3.6. BIG BEAR SOLAR OBSERVATORY

allows to observe simultaneously in three spectral lines (H\(\alpha\), H\(\beta\) and CaII 8542 Å) and slit-jaw H\(\alpha\) pictures on the videotape with a cadence of 25 images per second. Some parameters of the MFS are:

- The size of the grating: 90 × 100 mm\(^2\) with 600 lines/mm.
- The linear dispersion: 1 Å/mm in the second right order.
- The working spectral lines range covers the Balmer series, some helium lines, H and K lines and some important multiplets of iron and other metals:
  - H\(\alpha\): 6503 - 6623 Å.
  - D lines of Sodium and Helium: 5829 - 5949 Å.
  - H\(\beta\): 4797 - 4917 Å.
  - H\(\gamma\), FeI, FeII, TiI Multiplets: 4277 - 4397 Å.
  - H\(\varepsilon\), H\(\zeta\) and H and K lines of CaII, HeI: 3870 - 3990 Å.
  - Higher Balmer lines: 3640 - 3814 Å.

![Figure 3.9: The optical layout of the Multichannel Flare Spectrograph.](http://www.asu.cas.cz/~sos/mfs.html)

The spatial resolution is \(\leq 1\) arcsec.

3.6 Big Bear Solar Observatory

The Big Bear Solar Observatory (BBSO) is located in Big Bear Lake, California high in the San Bernardino Mountains. Being surrounded by the lake, the place has a stable atmosphere, which is essential for the solar observations.

The telescope array consists of three telescopes shown in Fig. 3.10: 10 inch refractor, 26 inch reflector and a 8 inch refractor. Each telescope has three optical benches for experiments. (See [http://bbso.njit.edu/Research/Halpha/ha_inst.html](http://bbso.njit.edu/Research/Halpha/ha_inst.html))

3.6.1 26 inches Telescope

The 26 inches telescope is a reflector vacuum solar telescope with a 66 cm main mirror, based on a typical Gregory Coudé design.

A simplified sketch of the optical layout is shown in Fig. 3.11 the focal length of the primary paraboloidal mirror (TM1) is 2.5 m, resulting in an f-ratio of f/3.85. The foci of the secondary elliptical mirror (TM3) are 0.25 and 3.25 m, producing a
CHAPTER 3. INSTRUMENTATION

(a) Front view of the telescope

(b) Back view of the telescope

Figure 3.10: (a): Front end of the BBSO telescope array showing 10 inch and 26 inch telescopes. The 8 inch refractor is hidden behind the 10 inch in this picture; (b): Back end of the 10 inch and 26 inch telescopes.

A magnification of 13, changing the system speed from f/3.85 to f/50 and providing a plate scale of 6.3” mm⁻¹ in the Coudé focus. The focal length of the elliptical secondary mirror is 0.23 m, leading to a 59.9 mm diameter pupil image, which is formed close to the folding mirror (TM3) in the Hat section (see Fig. 3.11).

Figure 3.11: Optical layout of the BBSO 65 cm vacuum reflector. Figure taken from [Denker et al. 2007].

In order to study the morphology of the active region NOAA 8471 (see Section 4.6), I used high resolution Hα and the I, U, Q and V Stokes parameters data from the OSL (a camera from the Orbiting Solar Laboratory project) and VMG (videomagnetograms) instruments respectively (See Table 3.5).

<table>
<thead>
<tr>
<th>Instrument</th>
<th>Data</th>
<th>Spatial Resolution</th>
</tr>
</thead>
<tbody>
<tr>
<td>OSL</td>
<td>High resolution Hα</td>
<td>1.16 arcsec/pixel</td>
</tr>
<tr>
<td>VMG</td>
<td>Stokes parameters</td>
<td>0.68 arcsec/pixel</td>
</tr>
</tbody>
</table>

Table 3.5: Instruments and data taken with the 26 inches telescope from BBSO.

3.6.2 10 inches Telescope

The 10 inches telescope is a refractor solar telescope (Fig. 3.12) with a 25 cm main mirror, built in 1970 at the Big Bear Observatory and removed in 2006. This telescope is being replaced by the New Solar Telescope (NST) at Big Bear Observatory (http://www.bbso.njit.edu/newtelescope/).

Usually it takes Hα full disk images at the same time than the 26 inches telescope. In Table 3.6 there are the characteristics of the data used in this thesis.
Table 3.6: Characteristics of the data taken with the 10 inches telescope from BBSO.

<table>
<thead>
<tr>
<th>Data</th>
<th>Spatial Resolution</th>
</tr>
</thead>
<tbody>
<tr>
<td>Full disk H$\alpha$</td>
<td>1.05 arcsec/pixel</td>
</tr>
</tbody>
</table>

Figure 3.12: Main observing floor and back end of the 10 inch telescope at BBSO. Under the white box near the middle of the picture is a videomagnetograph system. The optical circuit of the Multichannel Flare Spectrograph.

3.7 Kanzelhöhe Observatory

The Kanzelhöhe Observatory (Fig. 3.13(a)), in Austria, has a refractor telescope with a diameter of the entrance pupil of 100 mm and an effective focus length $f$ of 2000 mm. To study the Sun in H$\alpha$, it uses a band-pass filter at 6563 Å and a Zeiss Lyot filter with a FWHM of 0.7 Å as shown in Fig. 3.13(b). (See http://www.kso.ac.at/instrumente/instrumente_en.php).

I used the full disk H$\alpha$ images taken on 03 July 2002 in order to study the evolution of the active region NOAA 10019 (see Chapter 5). The images have a spatial resolution of 2.23 arcsec/pixel.

3.8 Nancay Radioheliograph Observatory

The Nancay Radioheliograph (NRH) consists of 44 antennas, situated at 2° east and 47° north (see Fig. 3.14). It allows fast 2D imaging from 150 to 450 MHz, but with a very low spatial resolution, 6 arcsec/pixel (Kerdraon & Delouis 1997).
3.9 Nobeyama Radio Observatory

The Nobeyama Radio Observatory (NRO) is a division of the National Astronomical Observatory of Japan (NAOJ). NRO is composed by different radio telescopes like:

- The 45 m radio telescope,
- The Nobeyama Millimeter Array (NMA),
- ASTE Submillimeter Telescope at Chile,
- The Nobeyama Radioheliograph
- The Nobeyama Radio Polarimeters

In Sections 3.9.1 and 3.9.2 I will detail the Nobeyama Radioheliograph (NoRH) and Radio Polarimeters (NoRP) respectively.

3.9.1 Nobeyama Radioheliograph (NoRH)

The Nobeyama Radioheliograph (NoRH) is a ground radio solar interferometer situated in Japan [Nakajima et al. 1994]. It consists of 84 parabolic antennas with 80 cm

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Figure 3.14: Nancay Radioheliograph: 44 parabolic antennas imaging the full Sun between 150 MHz and 450 MHz.

Figure 3.15: Nobeyama Radioheliograph: 84 parabolic antennas with 80 cm diameter.
3.9. NOBEYAMA RADIO OBSERVATORY

3.9.1 Diameter, separated 490 m in the east/west direction and 220 m in the north/south (Fig. 3.15). The first observation was in April 1992 and from June 1992 it takes 8 hours observation daily.

The data are full disk images of flux density (brightness) at 17 and 34 GHz with a spatial resolution of 10 and 5 arcsec/pixel respectively and a temporal resolution of 0.1 seconds during flare events or 1 second in steady mode.

3.9.2 Nobeyama Radio Polarimeters (NoRP)

The Nobeyama Polarimeters (NoRP, Fig. 3.16) started to work at 17 GHz in 1978; in 1983 it was added the 35 GHz frequency and the 80 GHz in 1984. The other frequencies (1, 2, 3.75, and 9.4 GHz) were working at Toyokawa observatory since 1979 and they were moved to Nobeyama in April 1994.

The polarimeters observe the full disk in multiple radio frequencies (1, 2, 3.75, 9.4, 17, 35 and 80 GHz), with a temporal resolution of 0.1 seconds. It records the flux and the polarisation at each frequency (Torii et al. 1979). NoRP has no imaging capabilities but instead provides the spatially integrated radio spectrum.

Figure 3.16: Nobeyama Radio Polarimeters.
OBSERVATIONS
Chapter 4

Flares in Lyα

4.1 Historical Overview

The Lyα line is the resulting emission of photons from the hydrogen atom when the electron falls from the energy level \( n = 2 \) to \( n = 1 \). It is the main line of the Lyman series and it is situated at 1215.67 Å, at the ultraviolet range of the spectrum.

Because the Lyα emission is a natural tracer of solar activity in the chromosphere and transition region, Lyα images can be used to study the morphology and evolution of active regions on the Sun, as well as the variability of solar irradiance.

The Lyα line has been of interest in flares as a possible diagnostic of proton beams. Orrall & Zirker (1976) proposed that charge exchange between beam protons and chromospheric hydrogen, and emission from the moving hydrogen atoms that result, would lead to enhanced emission in the wings of the chromospheric lines. More refined calculations (e.g. Canfield & Chang 1985; Fang et al. 1995; Zhao et al. 1998) suggested that this should be an observable effect. Early Skylab observations of the Lyα line profile by Canfield & van Hoosier (1980) detected a slight red asymmetry in the early phase of a flare, however the profiles were not in agreement with models of Canfield & Chang (1985), and the SOLSTICE observations of Brekke et al. (1996) showed a stronger Lyα blue wing. (A related search for alpha-particle beams using He II Lyα emission by Brosius (2001) also had a null result). These discrepancies could be explained in terms of various inclinations of the proton beam with respect to the magnetic field and the line-of-sight (Zhao et al. 1998). Henoux et al. (1995) and Zhao et al. (1998) conclude from their theoretical studies that non-thermal Lyα emission due to electron or proton beams could be detected best at the very early stages of the flare, when the coronal column mass is small. In later stages the beams have lost most of their energies when reaching the layers where Lyα is formed. Of course this is specific to Lyα and the effect of beams can be stronger on other lines (e.g. Lyβ).

On this basis, it seems plausible that line shifts and wing enhancements in the Lyα line profiles observed at later stages are due to bulk velocities of hydrogen atoms in the flaring atmosphere. There are few examples of theoretical studies incorporating both the effects of a non-thermal electron beam and of flows on the shape of the Lyα line profile. We refer the reader to the radiative hydrodynamic models of Allred et al. (2005) and our recent study in Chapter ??.

\(^1\)All the transitions associated to the hydrogen atom, between the orbital \( n \geq 2 \) to the orbital \( n = 1 \) are called the Lyman series. The transitions are named sequentially by Greek letters: Lyα is the resulting from \( n = 2 \) to \( n = 1 \), Lyβ, from \( n = 3 \) to \( n = 1 \), etc.
CHAPTER 4. FLARES IN $\text{Ly}\alpha$

A small number of flares observed in the $\text{Ly}\alpha$ passband early in the TRACE mission have been reported in the literature. For example Schrijver et al. (1999) and Schrijver (2001) discuss the downflows observed in $\text{Ly}\alpha$ during the late phase of a flare, and Fletcher et al. (2001) examine 1216 Å ejecta and chromospheric source movements at low cadence, just after the impulsive phase. Warren & Warshall (2001) mention that $\text{Ly}\alpha$ pre-flare brightenings are not co-spatial with subsequent flare hard X-ray emission sites. However, to our knowledge no detailed analysis of TRACE $\text{Ly}\alpha$ channel observations spanning the impulsive phase have been presented. We will explore the information available from the TRACE 1216 Å channel alone, and in combination with the TRACE 1600 Å channel, estimating the flare $\text{Ly}\alpha$ intensity for comparison with previous work.

4.2 TRACE UV “Contamination”

As shown in Figure 4.1, the TRACE 1216 Å channel has a broad spectral response, having approximately the same sensitivity in the $\sim$1500-1600 Å range as it does at $\sim$ 1216 Å (see Handy et al. 1999b). This unusual spectral response results from the convolution of a narrowband UV coating on the primary mirror at 1550 Å and a filter near the focal plane centered at 1216 Å. In the quiet sun, much of the contribution to the 1216 Å channel is continuum due to the double peak seen in Fig. 4.1, at the spectral response (see Fig. 4.1): one is located at $\sim$1216 Å and the second at $\sim$1550 Å. The 1500-1600 Å range, for which the 1600 Å channel is optimised, includes the strong C IV pair, as well as lines of Si II and continuum. It has essentially no response at wavelengths less than 1400 Å.

As we can see in Fig. 4.1, around 60% of the response in the 1216 Å channel is from the $\text{Ly}\alpha$ 1216 Å emission line while the balance is from UV emission near 1550 Å and longer wavelengths. Handy et al. (1999b) provide a simple method to remove this contamination by using a large number of solar irradiance data samples from Solar-Stellar Irradiance Comparison Experiment (SOLSTICE; Rottman et al. 1993), and by comparing TRACE 1216 Å images with $\text{Ly}\alpha$ images from the Very-
4.3. FLARES OBSERVED WITH THE TRACE 1216 Å CHANNEL

high resolution Advanced ULtraviolet Telescope (VAULT; Korendyke et al. 2001) sounding rocket flight of May 7, 1998. Their result is that pure Lyα intensity can be obtained by a linear combination of TRACE 1216 Å and 1600 Å channels:

\[ I_{\text{Ly} \alpha} = A \times I_{1216} + B \times I_{1600}, \]  

(4.1)

where \( I_{\text{Ly} \alpha} \) is the intensity of the ‘pure’ (i.e. corrected) Lyα emission line, \( I_{1216} \) and \( I_{1600} \) are the intensities as observed with TRACE 1216 Å and 1600 Å channels, respectively, and \( A \) and \( B \) are the fitting parameters. Coefficient \( A \) is expected to be near unity, while coefficient \( B \) is expected to be negative because \( I_{1216} \) is contaminated by the UV continuum, which can be well represented by \( I_{1600} \).

To obtain an initial guess for \( A \) and \( B \), Handy et al. (1999b) first used the SOLSTICE database of solar spectra covering a wide range of solar activity. They convolved a large set of SOLSTICE spectra with TRACE 1216 Å and 1600 Å channel response functions to obtain synthetic \( I_{1216} \) and \( I_{1600} \) values, and integrated the same set of spectra along the Lyα emission line profile to obtain Lyα values. A least-squares regression was then applied to these intensities to find best-fit coefficients of \( A = 0.97 \) and \( B = -0.14 \).

To further improve the coefficients, they then compared TRACE observations with a VAULT Lyα image. For \( I_{1216} \) and \( I_{1600} \) values in equation 4.1, they used the pixel intensities of TRACE 1216 Å and 1600 Å images, while for Lyα values, they used the pixel intensities of a VAULT Lyα image assuming that the VAULT Lyα image is not contaminated by non-Lyα emission (the Lyα filter of the VAULT has a bandpass of 150 Å). Fixing the coefficient \( A \) to be the value obtained from the SOLSTICE spectra, 0.97, they found that \( B = -0.105 \) results in the highest Spearman’s rank correlation \( r_s \) (Press et al. 1992) between both sides of eq. 4.1.

Kim et al. (2006) found the best-fit coefficients of \( A = 0.97 \) and \( B = -0.12 \), improving the results of Handy et al. (1999b), where \( B = -0.14 \), comparing the Lyα images with some observations from the Solar Ultraviolet Measurements of Emitted Radiation (SUMER; Wilhelm et al. 1995) instrument onboard Solar and Heliospheric Observatory (SOHO).

These coefficients apply to quiet Sun observations, and it is not clear that they can be applied to flare observations, in which the ratio of contributions in each of the TRACE channels may be different from their quiet Sun values, depending on the line and continuum excitations. The SOLSTICE flare observations of Brekke et al. (1996) showed a proportionally much larger increase in C IV emission than in Lyα. The continuum in both regions increased by about a factor 2. However, as noted above, because of the direction of the wavelength scan, the Lyα intensity was measured by Brekke et al. during the decay phase, whereas the C IV intensity was measured in the impulsive phase. In the absence of any further information we will use the Kim et al. (2006) prescription.

4.3 Flares Observed with the TRACE 1216 Å Channel

We searched in the TRACE flare catalogue for all the flares available in the TRACE 1216 Å and 1600 Å channels, in order to study the Lyα emission during flares. We found that there are few data available during flares (see Table 4.1) and that there is reasonable coverage in 1216 Å and 1600 Å for just one flare observed on 8 September 1999, and from the beginning till the maximum of another flare on 28 February 1999.

The 8 September 1999 flare was observed with an image size of 768 × 768 pixels; 1216 and 1600 Å images were interleaved with a cadence of approximately 5 seconds
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CHAPTER 4. FLARES IN LYα

per image at the very beginning and end of the event; there are also some white light, 171 and 195 Å images.

The 28 February 1999 flare was observed with an image size of 1024 × 1024 pixels; 1216 and 1600 Å images were interleaved with a cadence of approximately 5 seconds per image and there are also some 171 and 195 Å images.

4.3.1 The 8 September 1999 Flare

The M1.4 flare which occurred in active region NOAA 8690 on 8 September 1999 at 12:13 UT was observed by TRACE in different channels. Table 4.2 shows in detail the available TRACE data.

This flare was observed also by the Yohkoh satellite. The data available are:

- for SXT, every ≈ 20 sec from 12:13:41 to 12:14:07 UT (impulsive phase) and from 12:23:39 (gradual phase) till the end of the flare.

- for HXT, there are just 3 sets from 12:13:33 to 12:14:23 UT (impulsive phase).

We could not use the SOHO/MDI full disk magnetograms because the flare is situated near the limb of the Sun: SOHO/MDI measures the value of the component of the magnetic field along the line of sight; looking near the center of the Sun, this value is similar to the longitudinal component of the magnetic field, but near the limb of the Sun, this component is a mix of the longitudinal and transverse component, which is not possible to separate.

There are also some white-light and Hα images from the Big Bear Solar Observatory taken at 16:57:28 and 16:26:47 UT respectively (after the flare); besides, data of the previous day and the day after the flare are available.

4.3.2 The 28 February 1999 Flare

The TRACE data available for the M6.6 flare which occurred in active region NOAA 8471 on 28 February 1999 at 16:31 UT are listed in table 4.3. There are data in the TRACE 1216 Å and 1600 Å wavelengths till the maximum of the flare (pre-flare and impulsive phase), while there are no data for the gradual phase of the flare.

There are no Yohkoh data during the flare because at this time the satellite was eclipsed by the Earth.

Full disk MDI/SOHO magnetograms at Ni I 6767.8 Å, are available every 96 minutes from the day before to the day after the flare; the nearest magnetograms to the maximum of the flare are at 16:03 UT, before the flare and at 17:36 UT, after the flare. The MDI/SOHO full disk magnetograms are 1024 × 1024 pixels image with a plate scale of 1.98 arcsec/pixel.

We also have images from the Big Bear Solar Observatory:

- Hα full disk images acquired with the 25 cm telescope at the Big Bear Solar Observatory during the time interval 15:31:04 - 18:34:24 UT and 22:09:13 - 23:51:29 UT. The plate scale of the images is of 1.05 arcsec/pixel. The images were taken using a Hα filter, centered at the Hα line center (6562.8 Å) and using a bandpass of 0.25 Å.

- Hα high resolution (512 × 512 pixels) images acquired by the 65 cm telescope at the Big Bear Solar Observatory, with a plate scale of 1.158 arcsec/pixel. Further images are available for the previous day and the day after the flare.
### 4.3. FLARES OBSERVED WITH THE TRACE 1216 Å CHANNEL

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Table 4.1: List of flares observed with the TRACE 1216 Å channel.
Filter | Time Range (UT)
---|---
1216 | 07:52:13 - 15:54:12
1600 | 07:52:07 - 15:54:05
171 | 07:51:33 - 10:59:38
 | 11:46:17
 | 12:28:26
195 | 07:52:01 - 11:00:06
 | 11:46:44
 | 12:28:45
 | 13:14:46 - 15:50:00
white-light | 07:50:51 - 12:01:22
 | 12:01:22
 | 12:28:05

Table 4.2: TRACE data available for the flare occurred on 8 September 1999. The detailed time not ranged means that there is just one image at this time.

Filter | Time Range (UT)
---|---
1216 | 12:01:38 - 16:41:15
1600 | 15:00:54 - 16:41:08
171 | 16:10:54 - 16:41:32
195 | 13:00:09 - 16:41:42

Table 4.3: TRACE data available for the flare occurred on 28 February 1999.

(See Table 4.4). All the images were rotated by the P-angle. The images were taken using a Hα filter, centered at the center of the Hα line (6562.8 Å) and with a bandpass of 0.25 Å. For the 27 February observations, there are also images acquired along the Hα profile at ±0.25, ±0.5, ±0.75 and ±1 Å.

Day | Time Range (UT)
---|---
01-03-99 | 18:17:07 - 18:19:05

Table 4.4: BBSO Hα high resolution data available for active region NOAA 8471 from 27 February to 01 March 1999.

- High resolution (512 × 481 pixels) images of the 4 components of the Stokes parameters at 6103 Å acquired with the 25 cm telescope at the Big Bear Solar Observatory, with a plate scale of 0.683 arcsec/pixel. There are also some data available on the 27 February and 01 March 1999 (See Table 4.5).

The Stokes-I component is interpreted as a white light image, which gives us information about the morphology of the active region in the photosphere;

The P-angle is the angle between the geocentric north pole and the solar rotational north pole measured eastward from geocentric north (http://www.qsl.net/f5ru/glossary.htm). It ranges between ± 26.31 degrees.
4.4 DATA ANALYSIS

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<td>V</td>
<td>17:30:08 - 01:05:14</td>
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Table 4.5: High resolution BBSO Stokes I, Q, U and V data available for active region NOAA 8471 from 27 February to 01 March 1999.

from the Stokes-V component we can obtain the line-of-sight magnetogram [Landi Degl’Innocenti 1992].

4.4 Data Analysis

As a first step, it was necessary to correct the data for instrumental effects and to solve some pointing problems when comparing images from different satellites or even onboard the same satellite (e.g. images of the 1216 Å and the 1600 Å channels of TRACE which have an offset and slightly different plate scale as the focal position changes between wavelengths). The data was corrected also for the solar rotation.

4.4.1 TRACE

The original images taken by the TRACE telescope were corrected for instrumental effects (subtract the dark pedestal, correcting for exposure time, for radiation spikes or for saturated pixels) using the standard procedures included in the Solar Software (SSW; Freeland & Handy 1998) IDL routines provided by the TRACE team:

- Subtraction of the dark pedestal (Analog-Digital-Converter ADC) to an image: the pedestal and dark current value of the images is automatically obtained by the IDL routine TRACE_PREP in DN, where 1 DN is the basic data number unit which corresponds to an amplifier gain of 12 electrons/DN, and each electron corresponds to a detected photon that passes in the lumigen coating.

- Flat-field correction: it is a technique used to improve quality in digital imaging, removing artifacts from 2-D images that are caused by variations in the pixel-to-pixel sensitivity of the detector. The flat-field image is an uniformly illuminated image, with a typically quite short exposure time. Usually the telescope points to the quiet sun region. The flat-field correction is performed using the IDL routine TRACE_PREP.

- Correction for exposure time: not all images have the same exposure time, so we have to normalize them, dividing the counts of each pixel by the exposure time.
CHAPTER 4. FLARES IN LYα

Figure 4.2: There is a shift between 1216 and 1600 Å TRACE images taken on 08 Sept. 1999 at 12:00:20 UT: -2.7 pixels on x and 1.24 pixels on y because of changes in the focus when the telescope changes the filter (compare images (a) and (b)).

- Image normalisation to DN per sec: once we have corrected the images for the exposure time, the value of each pixel is given in counts/sec/pixel. If we are interested in measuring the intensity, we should normalize the region of the image to DN/sec by dividing the total amount of DN into a region by the size in pixels of that region. See for details Section 4.5.3.

- Alignment and pointing correction between two UV channels: an alignment and pointing correction between two UV channels is made under the assumption that the trace pointing does not vary much by using the /wave2point_correct keyword in TRACE_PREP, based on the wavelength dependent alignment corrections.

Because there is an offset between the TRACE images at 1216 Å and 1600 Å, we tried to correct it measuring the displacement of the network bright points between both images at the same time. In Fig. 4.2(a) we can see that the contours of the 1600 Å image are not completely overpositioned with respect to the bright points of the 1216 Å image. Instead in Fig. 4.2(b), where this displacement due to the offset between both channels is already corrected, the contours correctly overlay the darker regions.

4.4.2 Yohkoh

For the 8 September 1999 observations, Yohkoh entered in flare mode at 12:13:36 UT. The flare response provided full-resolution SXT images with 2.56" pixels of a region 2.6 × 2.6 arcmin² at a cadence of about 2 seconds. High-cadence HXT data were also taken. Unfortunately, coverage stops at 12:14:30, shortly after the flare mode onset, due to a data downlink. However there are sufficient data to reconstruct three images.

In order to compare TRACE images with Yohkoh data, we made a TRACE pointing correction, based on co-aligning the strongest TRACE and Yohkoh HXT sources, assumed to be flare footpoints. We estimated that this can be done with an accuracy of 5". The SXT pointing with respect to HXT is accurate to within 1".
4.4.3 SOHO/MDI

The data already calibrated and corrected for instrumental effects are available online at [http://soi.stanford.edu/production/time_range.html](http://soi.stanford.edu/production/time_range.html).

Comparing the MDI magnetograms relevant to the event occurred on 28 February 1999 with the images from TRACE or BBSO, it is also necessary to co-align the magnetograms with TRACE and BBSO images; for that we assumed that the pointing correction was the same for the magnetograms and the white light images from SOHO/MDI. We compared the SOHO/MDI white light image with the TRACE 1600 Å ones, in which sunspots were visible and we estimated the offset value between both images.

As already mentioned, for the 08 September 1999 event, as the flare is situated near the limb of the Sun, the MDI magnetograms cannot be used.

4.4.4 BBSO

We deduced the pixel size and the field of view of the data provided by BBSO using the following method: we compared the high resolution images of the Stokes V component with the full disk MDI/SOHO magnetograms by overlapping the contours corresponding to the penumbra edges of the main sunspots of the active region. From the best fit of the overlapping contours we determined the pixel size of all the images of the 4 components of the Stokes parameters. Then we compared the images of the Stokes I component with the high resolution Hα images taken at + 1.0 Å from the center of the line. In this case the best fit of the contours of the umbra and penumbra of the main sunspots provided us the pixel scale of images taken by the 65 cm BBSO telescope. Finally, by the comparison of the high resolution and the full disk Hα images we deduced the pixel scale for the Hα images taken by the 25 cm telescope.

4.5 Overview of the 8 September 1999 Event

4.5.1 Morphology of the Event

Active region NOAA 8690 appeared on 5 September 1999 as a small beta region on the north east of the limb of the Sun with arch type filaments, bright Hα plage and small sunspots. Fig. 4.3(a) shows its configuration in white light before the M1.4 flare, which is a sunspot group characterized by a βγδ magnetic configuration (from the BBSO Solar Activity Report). TRACE images in the 171 Å and 195 Å wavelengths from just before the flare show two small filaments (Figs. 4.3 (c) and (d)). The eastern most filament is visible also in emission in the Lyα channel (Fig. 4.3(b)), unlike the western filament, which is not clear either in emission or absorption.

The event starts around 12:05 UT with bright ejecta seen in TRACE 1216 Å. A pair of TRACE ribbons starts brightening at approximately 12:13:07 UT, with a maximum at 12:14:23 UT. The ribbons visible in Fig. 4.4 spread apart, and the ejecta remain bright until around 12:16:30 UT, after which they become faint and diffuse, and difficult to distinguish from the background plage (see difference images in Fig. 4.13). Between the brightest ejecta, seen in the left and middle columns of Fig. 4.4 and the footpoints there appears to be a void with a well-defined upper boundary, implying the ejection of a discrete structure rather than a continuous outflow. The cloud of ejecta, most clearly visible to the south and east of the ribbons, does not appear like expanding loops or a confined jet, or a coherent dense filament. It has the
appearance of a spreading semi-circular sheet of material with embedded “blobs”. This type of morphology is also seen in around 25% of flares observed in soft X-rays, and was classified as a spray by Kim et al. (2004). In the final phase of the flare, TRACE 1216 Å images show two ribbons slowly separating.

Fig. 4.4 shows the overlay of soft X-ray emission on 1216 Å images. The first two images in each row are from the impulsive phase, at 12:13:42 UT and 12:14:02 UT, and the third is from the gradual phase. The bright emission seen in the TRACE 1216 Å channel is spatially well correlated with the strongest emission seen in the SXT images. As the main emission in TRACE appears to come from footpoints, this is suggestive of early phase soft X-ray footpoints (as observed previously by
4.5. FLARE OCCURRED ON 08 SEPTEMBER 1999

Figure 4.4: Temporal evolution of the Yohkoh/SXT source (green contours), superimposed on the nearest TRACE 1216 Å images. The upper panels show Be 119 filter contours, and the lower panels show Al 112 filter contours. The contour levels are 1%, 5%, 10% and 50% of the peak counts in each SXT image.

Also apparent from these images is a faint SXR emission to the south-east of the strong footpoints, and co-spatial with the brightest part of the ejecta. This might imply that both hot and cool plasma are present in the ejecta. Later in the event the brightest SXR emission is co-spatial with the eastern ribbon. However, its position is also consistent with what would be expected of emission from the top of an arcade of hot flare loops joining the two sets of footpoints, as is usually found in the later phase of flares.

Figure 4.5: The 14-93 keV HXR source (blue contours), superimposed on the nearest TRACE 1216 Å (upper) and 1600 Å (lower) images.
Fig. 4.5 shows the hard X-ray flare sources overlaid on 1216 Å (upper panel) and 1600 Å (lower panel) TRACE emission. The HXR images are produced using the MEM algorithm in the Yohkoh/HXT software, integrating HXR counts from 14-93 keV. The HXR and TRACE 1216 Å footpoints again coincide well. Fig. 4.6 also shows an overlay of HXR emission on the nearest Ly$\alpha$ images, this time corrected for the UV contamination using Eq. (4.1). The use of the corrected images allows us to better visualise the correspondence between HXR source and the flare ribbons seen in Ly$\alpha$.

Fig. 4.7 shows the evolution of the flare at 171 Å. A few hours after the flare, a post-flare arcade is visible at coronal temperatures. As one would expect, the post-flare loops bridge the ribbons that are observed at 1216 and 1600 Å, expanding into and beyond the ‘void’ region that appeared below the UV ejecta, even taking the same shape as this void - compare loops in Fig. 4.7 left panel with void in Fig. 4.6 top right.

Figure 4.6: Temporal evolution of the 14-93 keV source (blue contours), superimposed on the nearest Ly$\alpha$ images corrected of ultraviolet contamination, using the equation 4.1.

### 4.5.2 Correction of UV "Contamination" on TRACE 1216 Å

As mentioned in Section 4.2, Handy et al. (1999b) assumed that a pure Ly$\alpha$ intensity can be obtained by a linear combination of TRACE 1216 Å and 1600 Å channels.

We used their estimation for the correction of the UV contamination in the 1216 Å TRACE channel provided by Eq. 4.1 and as we can see in Fig. 4.8, the superposition of the soft X-ray isocontours on the corrected Ly$\alpha$ images seems to fit better in the plasma that erupted in the eastern part of the Sun than on the flare ribbons.

We can see the overlay of the three different soft X-ray filters in Fig. 4.8 (Be 119 mm on the top, Al 112 mm in the middle and AlMg on the bottom) and of hard X-ray (the four channels) in Fig 4.6. Both images show that in the beginning the main X-ray source is well overlayed with the two ribbons, but then it is moving to the east of the Sun (compare Figs. 4.6(a) and (c)). This probably occurs because the X-ray source is situated in an upper layer, while with TRACE we are looking at the chromosphere, in a lower layer.

### 4.5.3 Estimation of the Ly$\alpha$ Intensity at the Flare Footpoints

It is possible to measure, in the 1216 Å TRACE flare image, the intensity in DN and convert it to erg s$^{-1}$, in order to compare our results with the theoretical values.
4.5. FLARE OCCURRED ON 08 SEPTEMBER 1999

(a) 171 Å at 15:33 UT
(b) 171 Å at 18:22 UT

Figure 4.7: 171 Å TRACE images from later on in the event.

Figure 4.8: Temporal evolution of the 0.25-4.0 keV source (green contours) for the three different filters: Be 119 (at the top), Al 112 (in the middle) and AlMg (at the bottom), superimposed on the nearest Lyα images corrected for ultraviolet contamination, using Eq. 4.1. The contour levels are the 2, 3, 10 and 50% of the peak counts for Be 119 mm, 0.2, 0.1 and 0.05% for Al 112 and 0.01, 0.05 and 0.08 for AlMg.
We selected a square around the flare region (between 300 and 550 pixels for x and between 200 and 450 for y) in a 1216 Å TRACE image at the impulsive phase of the flare (12:13:41 UT). Because in the TRACE image there are some saturated pixels, and because we want to exclude the intensity of the quiet sun, we excluded them before calibrating the image, in order to get the intensity associated to the footpoints of the flare; for that, we used an intensity range of: 4090 DN for the maximum and 1200 DN for the minimum.

As already mentioned, the original images taken by TRACE must be corrected for instrumental effects. We calibrated the images (without the saturated pixels) using the Standard Solar Software (SSW) provided by the TRACE team (Freeland & Handy 1998) (see Section 4.4).

After the calibration, we measured the intensity at the flare footpoints of the TRACE 1216 Å image (9.3 × 10⁵ DN); dividing this value by the exposure time of the image, we obtain the intensity in DN s⁻¹ (3.1 × 10⁶). Knowing the size of the square around the flare region, it is possible to obtain the intensity in DN/s/pixel (7748.5). Taking into account the response of the TRACE filter, we know that 1 photon s⁻¹ cm⁻² sr⁻¹ is 8.0 × 10⁻¹⁵ DN s⁻¹ pixel⁻¹, so our footpoints intensity is 9.7 × 10¹⁷ photon s⁻¹ cm⁻² sr⁻¹. Assuming that all the energy comes from Lyα photons and knowing that the energy of a Lyα photon is \( h\nu = 1.63 \times 10^{-11} \text{ erg} \), the intensity at the flare footpoints is around 2.68 × 10⁷ erg s⁻¹ cm⁻² sr⁻¹.

This value is an upper limit, because there is some contribution from the UV as well. We followed the same steps for the TRACE 1600 Å image at 12:13:37 UT and we recalculated the intensity at the Lyα image taking into account the assumption that there is a linear combination between both TRACE filters, so the intensity at the flare footpoints in the Lyα image corrected from UV contamination is 9.0 × 10⁵ DN/s, lower than the intensity in the TRACE 1216 Å image at 12:13:41 UT. Following the previous steps, the power is 2.4 × 10²⁵ erg s⁻¹. Table 4.6 shows the resulted values.

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Table 4.6: Intensity estimated at the flare footpoints of the TRACE 1216 Å and 1600 Å images, during the impulsive phase.

4.5.4 Light Curves

Fig. 4.9 shows the time profiles of the TRACE 1216 Å and 1600 Å channels, and the corrected Lyα, in counts per second. These were obtained by thresholding the normalised exposures at a level of 4000 counts per second, thus isolating the brightest flare sources. The footpoint sources were very intense, but saturation was mostly avoided as the TRACE automatic exposure control limited the exposure times to between 0.2 and 0.3 seconds in the 1216 Å channel, and between 0.04 and
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The TRACE 1216 Å and 1600 Å light curves show that the 1600 Å flux rises earlier and more rapidly than the 1216 Å flux, and also peaks ~1 minute earlier. The TRACE 1216 Å light curve decays more rapidly. During the flare observed by Brekke et al. (1996) the Lyα line remained stronger overall, although the C IV lines increased in intensity by a much larger factor than did the Lyα line. We can deduce that the C IV contribution to the 1216 Å channel during the flare is rather small, and the two light curves primarily reflect the behaviour of the Lyα and C IV lines. There is evidence for more early phase emission in C IV, with a somewhat more impulsive behaviour. This might indicate that the atmospheric layers responsible for producing the C IV emission are heated/excited before the Lyα emitting layers. The corrected Lyα also rises earlier than 1600 Å or 1216 Å. This Lyα time information must be treated with caution because we time-interpolated intensities in the two TRACE filters, and because the correction we applied might not correctly estimate the C IV contribution, and bias the corrected Lyα light curve.

The HXT light curves from the early impulsive phase and in the decay phase are shown in Figure 4.10. Only the earliest part of the impulsive phase was observed before the Yohkoh data downlink, followed by spacecraft night. The flare was in its decay phase when observations started again. The Yohkoh/SXT light curve (Fig. 4.11) demonstrates the usual steep rise in soft X-rays at the beginning of the flare, and then the slow decay as the spacecraft comes out of eclipse.

4.5.5 Flare Hard X-Ray Energetics

Yohkoh/HXT observed this flare only for the first minute of the impulsive phase. However this provides enough data to calculate the electron beam power under the assumption of a chromospheric collisional thick target. The four HXT energy channels L, M1, M2 and H showed strong impulsive bursts. We fit the photon spectrum with the Eq. 4.2:

\[ I(\epsilon) = A\epsilon^{-\gamma}\ \text{photons cm}^{-2}\ \text{s}^{-1}\ \text{keV}^{-1}, \]  

where \( I \) is the photon flux at the energy \( \epsilon \) in keV, \( \gamma \) is the power law index and \( A \) is a normalisation constant. It is then possible to deduce the instantaneous electron en-
energy flux at injection, above some “cut-off” energy $E_c$, assuming that the power-law part of the spectrum is generated by electron-proton bremsstrahlung in a collisional thick target process. Assuming the Bethe-Heitler cross section, the following can be used (Fletcher et al. 2007):

$$P(E \geq E_c) = 5.3 \times 10^{24} \gamma^2 (\gamma - 1) B\left(\frac{\gamma}{2} - \frac{3}{2}\right) AE_c^{1-\gamma},$$

where $B$ is the beta function and $A$ is the normalization constant of the photon spectrum $I(\epsilon)$.

In Fig. 4.12 we show the X-ray photon spectrum derived from the Yohkoh/HXT data at the maximum of the hard X-ray flare. The spectrum is well fitted by a power law with index 3.40 which indicates a very hard spectrum. This result suggests that the hard X-ray emission above 14 keV was produced by nonthermal electrons. We tried fitting both a power law plus thermal spectrum and a single thermal spectrum, but the single power-law provided a significantly better fit.

The total energy in the electron spectrum depends on the value of $E_c$ which cannot be determined from this spectrum. Instead we use three values of $E_c$ and calculate the instantaneous beam power carried by electrons above this energy. The spectrum implies that it might be reasonable to assume a low energy cutoff as low as 15 keV, though values of 20-25 keV are more typically assumed. In Table 4.7 we present the parameters derived from the single power law fit.

<table>
<thead>
<tr>
<th>Power law Normalization</th>
<th>$P \geq E_c = 15$</th>
<th>$P \geq E_c = 20$</th>
<th>$P \geq E_c = 25$</th>
</tr>
</thead>
<tbody>
<tr>
<td>index $\gamma$</td>
<td>constant $A$</td>
<td>(10$^{27}$ erg s$^{-1}$)</td>
<td>(10$^{27}$ erg s$^{-1}$)</td>
</tr>
<tr>
<td>3.399 ± 0.008</td>
<td>2.43 ± 0.08 x 10$^5$</td>
<td>8.61 ± 0.28</td>
<td>4.31 ± 0.14</td>
</tr>
</tbody>
</table>

Table 4.7: Parameters derived from Yohkoh/HXT data.

The values in Table 4.7 are reasonable for a small M-class flare such as the one studied here, and are also adequate to power the Lyα losses calculated in Section 4.5.3.

Figure 4.10: Time profile of corrected TRACE Lyα flux shown as pink triangles, and Hard X-Ray counts/subcollimator/second taken in the early impulsive phase with Yohkoh/HXT. The four different channels represented are 14-23 keV (L), 23-33 keV (M1), 33-53 keV (M2) and 53-93 keV (H). The inset shows the HXT time profiles at the beginning of the flare at higher time resolution.
Figure 4.11: Time profile of the Soft X-ray flux in counts/second taken in the early impulsive phase and at the end of the gradual phase with Yohkoh/SXT. The three different channels represented are Dagwood Sandwich (2.4-32 Å), Be 119 mm (2.3-10 Å) and Al 12 mm (2.4-13 Å).

Figure 4.12: Hard X-ray spectrum derived from the channel M2 of Yohkoh/HXT data at the maximum of the flare.
4.5.6 The Filament Ejection and Chromospheric Disturbance

The TRACE UV data allowed us to examine the evolution of extended flare region at very high spatial and temporal resolution. A series of difference images from the 1216 channel is shown in Fig. 4.13. The time of the reference image is 12:05:00 UT, the time at which the 1216 Å series began, and we have selected only a subset of the images to show the overall evolution. The flat-fielded and normalised TRACE images are first cross-correlated to correct for drift and solar rotation, and then the reference image is subtracted from following images, such that parts of the difference image appearing bright indicate enhanced emission. To the north west of the main flare brightening, at image pixel location ~ (360, 300) there is a dark region corresponding to the disappearance of some loops just visible in the first image. This could be due to the loops being ejected from the field of view within a minute or so (as no traveling bright feature is seen), losing mass in some other way, or being heated out of the 1216 Å passband. The ejecta appear as a ragged cloud expanding to the south east. This starts off bright and then fades in each subsequent image, presumably as the cloud expands. Around 12:14:41 UT and for a couple of minutes thereafter two patches at pixel location ~ (400, 220) and (400, 380) become bright.

In the three images following 12:15:31 the brightenings form a rough, diffuse semicircle, reminiscent of an ‘EIT wave’, but very much fainter. The cloud expanding to the south east is structured on a scale of ~ 20”, with brighter ‘knots’ within a more diffuse background. At 12:17:35, three or possibly four bright ‘knots’ of emission are apparent in the cloud, moving to the south east. These are resolved by TRACE, and have a dimension at that time of approximately 20”. The knots themselves also grow.

Figure 4.13: 1216 Å base difference images showing the large-scale evolution of the flare. The range shown from white to black is ±200 DN/s. The first image is the reference image used to perform the difference maps. In these images, north is at the top and west at the right.
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fainter and expand. A rough estimate of the projected speed of the east-most knot is $300 \pm 50 \text{ km} \cdot \text{s}^{-1}$. Eventually the cloud of emission is no longer detectable above the background. We interpret the expanding brightenings to the south-east as moving, emitting gas. The ejection to the south east is also faintly visible in soft X-rays at around 12:13 - 12:14 UT, so we conclude that the ejection contains both hot and cool material. However the brightenings to the west are more static (e.g. around pixel position (400,420)), and may include some component of network emission which has been enhanced by the passage of the disturbance.

Figure 4.14 shows the diffuse expanding emission in more detail. Contour levels of [60, 200] DN/s show the bright core and the diffuse, expanding outer portion of the ejecta. As previously, the 1216 Å emission in the knots can be estimated from the TRACE DN/s values: a value of 60 DN/s corresponds to an intensity of approximately $1.2 \times 10^6 \text{ erg cm}^{-2} \text{ s}^{-1}$.

The emission at Lyα wavelengths suggests cool plasma in the ejecta, but as noted in Section ??, the ejection to the south east is also faintly visible in soft X-rays at around 12:13 - 12:14 UT. Unfortunately the SXR spatial resolution does not allow us to compare in detail the spatial distribution of the 1216 Å and SXR emitting plasma, so we cannot say whether they are co-spatial or whether there are hotter and cooler clumps. It is possible that the ejecta are primarily hot, with a small fraction of neutral H atoms efficiently scattering the intense radiation around 1216 Å from the flare site.
4.5.7 Discussion of the Results Obtained from the Analysis of the 8 September 1999 Flare

The study of this event presents the first examination of a flare in TRACE 1216 Å at high spatial and temporal resolutions. We applied the empirical correction of Handy et al. (1999b), Kim et al. (2006) in an attempt to separate the Ly\(\alpha\) radiation from the UV contamination in the 1216 Å channel, and in doing so have shown that the majority of the Ly\(\alpha\) emission originates from the flare footpoints, which are compact and co-spatial with the HXR sources. Erupting material is also visible in the corrected Ly\(\alpha\) images, as well as in the 1216 Å channel. The corrected Ly\(\alpha\) peaks at approximately the same time as the HXR (within a minute or so) but unfortunately we do not have sufficient time coverage in HXT or time resolution in the TRACE images to say more than this. From a knowledge of the TRACE response function we have calculated the power radiated in the Ly\(\alpha\) footpoints, and shown that this can be provided by the electrons which produce the observed HXR radiation, under the assumption of a collisional thick target chromosphere.

The event was associated with the eruption of a small filament that was visible in the active region before the flare. During the eruption, the material emitted in soft X-rays as well as at the cooler TRACE wavelengths. The filament before the flare was presumably at temperatures of \(\sim 10,000\) K, so the SXR emission implies that the filament heated as it erupted. The resolution of the SXT images is not sufficient to say whether the SXR emission is uniform or clumpy. In TRACE, the erupting material looked sheet-like with some embedded brighter knots which spread and fade. The erupting material could be mixed blobs of hot and cool material, or conceivably mostly heated filament material but still with sufficient neutral hydrogen to scatter UV flare radiation from the chromosphere. This might be the case if the hydrogen in the erupting filament does not have time to reach ionisation equilibrium.

In summary, this study has confirmed much of what was previously known about flares in Ly\(\alpha\), including the typical Ly\(\alpha\) power radiated, and also demonstrated its close spatial and temporal association with the primary sites of energy release at the chromosphere. One of the most intriguing part of the observation is the filament eruption, which is already well underway before the flare footpoints brighten substantially, and which shows a morphology perhaps quite unlike what one would expect. It is neither rope-like - such as one might expect for the ‘core’ of a classical 3-part CME, or bubble-like, such as one might expect for the ‘front’ of a classical 3-part CME. In general, the flare events which are triggered by filament eruptions can be explained by the standard flare model CSHKP (Carmichael 1964b, Sturrock 1966, Hirayama 1974, Kopp & Pneuman 1976).

The fact that the 1216 Å channel and the corrected Ly\(\alpha\) emission both show the flare footpoints and the early stages of the onset of the eruption make this a promising channel for future observation of flares and eruptive events. Estimated Ly\(\alpha\) count rates are high from a small M flare such as this, which should enable high cadence imaging for similar or smaller events with future instruments.

4.6 Overview of the 28 February 1999 Event

The active region NOAA 8471 was observed by Big Bear Solar Observatory (BBSO) in white light (WL) and in the H\(\alpha\) line and by TRACE at different wavelengths (1216 Å, 1600 Å, 171 Å and 195 Å), which allowed the study of the morphology of the event, and by MDI on board SOHO, which allowed the study of the active
4.6. FLARE OCCURRED ON 28 FEBRUARY 1999

4.6.1 Morphology of the active region

The active region NOAA 8471 appeared on 23 February 1999 like a small beta region with arch type filaments, bright Hα increasing plage and small sunspots on the north east of the limb of the Sun. It was observed for 10 days and it had associated flare events (BBSO Solar Activity Report). On 28 February when an M6.6 flare occurred, the active region had coordinates N29 W12. The evolution of the active region, deduced by BBSO Solar Activity Report, is shown in Table 4.8.

In Fig. 4.15 we report a sequence of MDI magnetograms (top row) and TRACE 171 Å images (bottom row) to show the rapid evolution of the active region during the time interval 16:00 UT on 27 February to 16:00 UT on 28 February. In particular, we can see, comparing Fig. 4.15(a) with Fig. 4.15(b) that some knots of negative polarity in the central part of the active region appear to shift towards east, indicating the presence of shearing motions in this direction; comparing Fig. 4.15(d) with Fig. 4.15(e), we see that the morphology of the active region in the corona changed significantly during the same time interval.

The comparison of the MDI magnetograms acquired on 28 February at 01:35 UT and 16:03 UT (Figs. 4.15(b) and 4.15(c)) shows that the shearing motion is still going on in the central part of NOAA 8471, as well as the appearance of an extended region of negative polarity at position [150, 560] arcsec, which will give rise to a spot (see below). Comparing the TRACE 171 Å images acquired on 28 February at 00:06 UT and 16:03 UT (Figs. 4.15(e) and (f)) we can see that during this time interval some relatively small bright loops appeared in the central region (where the shearing motions took place), while the larger loops visible in Fig. 4.15(e) in the southern part of the active region are not anymore present in Fig. 4.15(f).

The white light image in Fig. 4.16(a), acquired at BBSO on 28 February 1999 at 17:49:52 UT, shows a sunspot group characterized in the western side by a main

<table>
<thead>
<tr>
<th>Day</th>
<th>Position</th>
<th>Magnetic Activity</th>
</tr>
</thead>
<tbody>
<tr>
<td>28 Feb. 1999</td>
<td>N28 W09</td>
<td>βδ class Region grow.</td>
</tr>
<tr>
<td>01 March 1999</td>
<td>N29 W25</td>
<td>βγ class Decaying phase. δ configuration no longer present.</td>
</tr>
<tr>
<td>02 March 1999</td>
<td>N29 W38</td>
<td>βγ class Slow decayed, but remained βγ magnetic class.</td>
</tr>
<tr>
<td>03 March 1999</td>
<td>N29 W52</td>
<td>βγ class Continued decay phase.</td>
</tr>
<tr>
<td>04 March 1999</td>
<td>N29 W65</td>
<td>βγ class Continued decay phase.</td>
</tr>
<tr>
<td>05 March 1999</td>
<td>N29 W85</td>
<td>βγ class Continued decay mode. Very close to limb.</td>
</tr>
</tbody>
</table>

Table 4.8: Evolution of the active region NOAA 8471.
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Figure 4.15: Top row: SOHO/MDI images acquired at (a): 16:03 UT on 27 Feb.; (b): 01:35 UT on 28 Feb.; (c): 16:03 UT on 28 Feb. The blue and green contours indicate the negative and positive magnetic field at ± 200, 1000 G. Bottom row: TRACE 171 Å images acquired at (d): 16:03 UT on 27 Feb.; (e): 00:07 UT on 28 Feb.; (f): 16:03 UT on 28 Feb. The images have a field of view of $1.81 \cdot 10^5 \times 1.09 \cdot 10^5$ km$^2$.

sunspot with a δ-configuration; some pores, with both positive and negative polarities, are present in the central and eastern side (compare with Fig. 4.16(c)).

The Hα image (Fig. 4.16(b)), taken at BBSO on 28 February 1999 at 17:49:48 UT, shows a bright facular region, and several filaments crossing the active region from the east to the west; the arrow remarks the presence of the filaments 1 and 2, which play an important role in the evolution of the flare. Similar inspection of images acquired at other wavelengths (i.e. TRACE 171 and 195 Å) indicates the presence of the corresponding EUV filament channels (arrows in Fig. 4.20 (c)-(d)).

In Fig. 4.16(c), the BBSO high resolution line-of-sight magnetogram, taken on 28 February 1999 at 17:49:52 UT, shows the magnetic configuration of the active

3 A sunspot has a δ configuration when both positive and negative polarities are inside the same penumbra

Figure 4.16: (a): White light image acquired at BBSO at 17:49:52 UT on 28 February 1999, where there is a sunspot with a δ-configuration; (b): High resolution Hα image acquired at BBSO at 17:49:48 UT, showing several filaments; (c): Magnetogram acquired at BBSO, showing the magnetic configuration of the active region at 17:49:52 UT. The images have a field of view of $4.74 \cdot 10^5 \times 1.12 \cdot 10^5$ km$^2$. 
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region, which appears quite simple, with a clear separation between the opposite polarities. Comparing both white light and Hα images with the magnetogram, we can see that the main sunspot, characterized by a $\delta$ configuration, is situated over the region encircled in Fig. 4.16(c), while the brightest facular region visible in the eastern part of the Hα image corresponds to the negative polarity of the active region.

In Figure 4.17 we can see a sequence of Hα images, where we can distinguish the bright facular pattern of NOAA 8471 and several filaments. On 27 February, at 18:59:24 UT (Fig. 4.17(a)), a main arched filament (indicated with arrow 1) and several smaller ones can be seen. On 28 February, at 17:51:18 UT (Fig. 4.17(b)), after the impulsive phase of the flare, the active region is very bright and the filament 1 and other smaller ones are still visible. The next day, on 01 March, at 18:18:05 UT (Fig. 4.17(c)), the filament 1 is still visible and there is evidence of other smaller filaments at the center and at the periphery of the active region.

4.6.2 Morphology of the event

In active region NOAA 8471, an M6.6 GOES class flare occurred on 28 February 1999, starting at 16:31:00 UT and reaching its peak at 16:38:35 UT. Fig. 4.18 shows the X-ray flux measured by the GOES satellite: after the flare peak there is an initial rapid decrease and later a much smoother gradual phase, lasting about 40 minutes.

A coronal mass ejection (CME) was observed by LASCO/SOHO C2 coronagraph at 17:54:05 UT. The corresponding time of the CME initiation calculated with the height-time measurements is 16:42 UT (see http://cdaw.gsfc.nasa.gov/CME_list/). The CME was associated with the M6.6 class flare in NOAA 8471.

Fig. 4.19 shows the time profiles of the TRACE 1216 Å, 1600 Å and corrected Ly$\alpha$ intensity, in counts per second. The 1600 Å intensity rises earlier and more rapidly than the 1216 Å intensity and it peaks $\sim$ 1.5 minutes later. During the decay phase, the intensity at 1216 Å decreases less rapidly than at 1600 Å. The corrected Ly$\alpha$ values are smaller than the 1600 Å and 1216 Å values. However, we must take into account that the Ly$\alpha$ emission reported in Fig. 4.19 must be treated with caution because we time-interpolated intensities in the two TRACE channels.

In Fig. 4.20 we can see the flare signatures in the early stage of the impulsive phase and at the peak in the 1600 Å, 1216 Å, 171 Å and 195 Å wavelengths, respectively.

In the early stage of the impulsive phase (at 16:33 UT), TRACE 1600 Å and 1216 Å images (Fig. 4.20(a) and (b)) show a sudden brightness increase at location [140:170; 550:560], while the 171 Å and 195 Å images (Figs. 4.20(c) and (d)) show bright areas also in the eastern side of the active region.
In Fig. 4.20(e), (f), (g) and (h) we can see the increased emission at the peak of the flare (16:41 UT) in different layers of the atmosphere: the TRACE 1600 Å (Fig. 4.20(e)) shows the flare configuration in the chromospheric layer, TRACE 1216 Å in the upper chromosphere (Fig. 4.20(f)), and TRACE 171 Å and 195 Å (Figs. 4.20(g) and (h)) are associated with the intensity emitted in the transition region and the corona.

At the peak of the impulsive phase of the flare, the bright emission is extended over a greater area [50:200; 520:580] at all wavelengths. The 1216 Å TRACE images show the active region at a deeper layer than the 171 Å, and we can see in Fig. 4.21 where we overplotted the 1216 Å contours at around 400 DN/s over the nearest 171 and 195 Å TRACE images that the flare
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Figure 4.20: 1600 Å, 1216 Å, 171 Å and 195 Å TRACE images at the beginning of the impulsive phase (a), (b), (c), (d) and at the maximum of the flare (e), (f), (g), (h). The field of view is $2.10 \cdot 10^5 \times 2.03 \cdot 10^5$ km$^2$.

region is well correlated at different atmospheric layers.

4.6.3 Flare evolution in the Hα line center

Fig. 4.22 reports the flare evolution in the chromosphere, in the center of the Hα line, using BBSO full disk images. Fig. 4.22(a) shows the active region about 8 minutes before the beginning of the impulsive phase. We can see that some filaments are present inside the bright facular region. If we compare this image with the BBSO

Figure 4.21: Overlay of the 1216 contours over the 171 and 195 Å TRACE images closer in time. The contours shown are 400 DN/s.
Figure 4.22: Flare evolution from 10 minutes before the beginning of the impulsive phase to 40 minutes after the flare in Hα, using a selected area of full disk BBSO images; the green arrows indicate the position of the two main filaments associated to the flare. The FOV of the images is $2.28 \times 10^5 \times 2.28 \times 10^5$ km².

magnetogram reported in Fig. 4.16 (c), we can see that these filaments outline the main polarity inversion line. In particular, we want to highlight again the presence of a filament at pixel location [170:195; 120:145] (filament 1) and another one at [100:140; 160:165] pixels (from now on called filament 2), indicated by arrows in Fig. 4.22 (a), because these features will have a key role in the flaring process.

In the following image (Fig. 4.22 (b), acquired at 16:35:49 UT), we can see that the site close to filament 1 becomes brighter, with a clear wave-like pattern. Later on the bright area increases its size and extends toward east, until it reaches the location where filament 2 is located. At 16:38:53 UT (corresponding to the time of
the peak of the flare according to the X-ray flux measured by GOES) in the Hα images, the two filaments are not visible anymore, as the brightness increase is now distributed along all the region located between [60:240, 120:175] (see Fig. 4.22(e)).

At approximately [100:150; 100:150] pixels, since 16:40:23 UT (Fig. 4.22(f)) an increasing bright area extends toward the south-east and, starting from 16:53:22 UT (Fig. 4.22(h)) the region of increased brightness at the site of filament 1 assumes more clearly the shape of two ribbons, that are generally observed on both sides of erupting filaments and that outline the footpoints of a magnetic arcade (compare Fig. 4.22(h) and Fig. 4.22(i)).

### 4.6.4 Flare evolution at 1216 Å and 1600 Å

A series of base difference images from the 1216 Å channel is shown in Fig. 4.23 (uncorrected for UV contamination coming from the TRACE 1600 Å emission), where the first image is the reference image used to perform the difference maps.

The time of the base image is 16:00 UT, ∼ 0.5 hour before the beginning of the flare. We have selected only a subset of the images to show the overall evolution. The flat-fielded and normalised TRACE images are first cross-correlated to correct the shift due to the solar rotation, and then the base image is subtracted from following images (see Section 4.4.1).

At 16:20 UT (Fig. 4.23(b)) we can see at the extreme western border of the active region ∼ [220; 590] arcsec, a very bright structure resembling a vertical upward jet of plasma. At the same time, at the position [140:170; 560:575] arcsec, we can see a brightness increase having a wave-like shape (compare with the same position in Figs. 4.20(a)-(d) and with Fig. 4.22(b)). In Fig. 4.23(c), the bright wave-like feature is still visible. This structure becomes more and more extended in the following images, moreover, in 1 minute and 15 seconds, (See Fig. 4.23(c) and (d)), an increase of brightness appears also in the region [79:106, 567:579], growing in time. We can see in Figs. 4.23(c) and (f) that later the bright area increases along the eastern part of the active region [60:112, 525:590], assuming a very diffuse pattern as time goes.

In the region centered at [225,590] the bright jet, firstly identified in Fig. 4.23(b), shows alternate phases of higher and lower emission, but it is always active till the maximum of the flare. Interestingly, this jet-like structure reaches greater length as time goes.

The analysis of TRACE images acquired at 1600 Å (Fig. 4.24), 171 Å (Fig. 4.25) and 195 Å shows a similar evolution of the flare signatures also at these wavelengths.

### 4.6.5 Estimation of the Lyα intensity at the flare footpoints

Following the same steps described in Section 4.5.2, we corrected the 1216 Å TRACE image from the UV contamination.

We selected a square around the flare region (from 98 to 223 arcsec for x and between 502 and 627 arcsec for y) in a 1216 Å TRACE image at the impulsive phase (16:38:45 UT), for a threshold range between 1200 and 4090 DN, to avoid the saturated pixels.

We calibrated the image following the same steps as for the other flare, as explained in Section 4.4.

After the calibration, we measured the 1216 Å intensity at the flare footpoints (1.3 × 10^7 DN): dividing this value by the exposure time of the image, we obtain 5.3 × 10^6 DN s^{-1}; knowing the size of the square around the flare region, the intensity is 996.4 DN s^{-1}pixel^{-1}. Taking into account the response of the TRACE filter, our
footpoints intensity is $1.2 \times 10^{17}$ photon s$^{-1}$ cm$^{-2}$sr$^{-1}$. Assuming that all the energy comes from Ly$\alpha$ photons, the intensity at the flare footpoints is around $2.0 \times 10^{16}$ erg s$^{-1}$cm$^{-2}$sr$^{-1}$. We calculated the intensity in erg s$^{-1}$sr$^{-1}$ ($2.1 \times 10^{21}$). Assuming isotropic emission, the power of the flare footpoints in Ly$\alpha$ is $2.7 \times 10^{22}$ erg s$^{-1}$. This value is the power from one pixel, then multiplying by the number of pixels in the selected square, we obtain $1.4 \times 10^{26}$ erg s$^{-1}$.

As for the previous flare, this value is an upper limit, because the TRACE 1216 Å channel is a broad channel which includes a UV ‘contamination’ as well. We estimated the UV ‘contamination’, calculating the intensity at the TRACE 1600 Å image acquired at 16:38:37 UT, taking into account the offset between the TRACE 1216 and 1600 Å images and following the previous steps.

Using Eq. 4.1, we recalculated the intensity at the flare footpoints in the Ly$\alpha$ image already corrected from UV: $2.9 \times 10^{6}$ DN s$^{-1}$, a value very close to the intensity in the TRACE 1216 Å image. Following the previous steps, the power is $8.1 \times 10^{25}$ erg s$^{-1}$, a factor of 0.6 smaller than the one measured in the 1216 Å image. Table 4.9 shows the resulted values.
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Figure 4.24: 1600 Å base difference images obtained subtracting the image acquired at 16:10:13 UT to the others. The range shown from white to black is ± 1200 DN/s. The FOV is 2.10 · 10^5 × 2.03 · 10^5 km^2.

<table>
<thead>
<tr>
<th>TRACE 1216 Å at 16:36:00 UT</th>
<th>TRACE 1600 Å at 16:35:53 UT</th>
<th>Lyα corrected at 16:36:00 UT</th>
</tr>
</thead>
<tbody>
<tr>
<td>DN s^-1</td>
<td>5.3 × 10^8</td>
<td>1.5 × 10^9</td>
</tr>
<tr>
<td>erg s^-1</td>
<td>1.4 × 10^{26}</td>
<td>4.2 × 10^{26}</td>
</tr>
</tbody>
</table>

Table 4.9: Intensity estimated at the flare footpoints of the TRACE 1216 Å and 1600 Å images, during the impulsive phase.

4.6.5.1 Comparison Between the Two Flares (08 September 1999 and 28 February 1999)

As shown in Table 4.10 for the M6.6 flare (28th February 1999), the power at the flare footpoints in Lyα greater than for the M1.4 flare (8th September 1999); this is consistent with the fact that this M6.6 class flare is more energetic in X-rays than the M1.4 class flare previously studied.
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Figure 4.25: 171 Å base difference images obtained subtracting the image acquired at 16:10:54 UT to the others. The range shown from white to black is ± 30 DN/s. The FOV is 2.10 \(\times\) 2.03 \(\times\) 10\(^5\) km\(^2\).

<table>
<thead>
<tr>
<th>Units</th>
<th>M6.6 Flare</th>
<th>M1.4 Flare</th>
</tr>
</thead>
<tbody>
<tr>
<td>DN s(^{-1})</td>
<td>2.9 (\times) 10(^6)</td>
<td>9.0 (\times) 10(^5)</td>
</tr>
<tr>
<td>DN s(^{-1})px(^{-1})</td>
<td>502.2</td>
<td>8890.06</td>
</tr>
<tr>
<td>ph s(^{-1})cm(^{-2})sr(^{-1})</td>
<td>6.2 (\times) 10(^{16})</td>
<td>1.1 (\times) 10(^{18})</td>
</tr>
<tr>
<td>erg s(^{-1})cm(^{-2})sr(^{-1})</td>
<td>1.0 (\times) 10(^6)</td>
<td>1.8 (\times) 10(^7)</td>
</tr>
<tr>
<td>erg s(^{-1})sr(^{-1})</td>
<td>1.1 (\times) 10(^{21})</td>
<td>1.9 (\times) 10(^{22})</td>
</tr>
<tr>
<td>erg s(^{-1})</td>
<td>1.3 (\times) 10(^{22})</td>
<td>2.4 (\times) 10(^{23})</td>
</tr>
<tr>
<td>total power</td>
<td>8.1 (\times) 10(^{25})</td>
<td>2.4 (\times) 10(^{25})</td>
</tr>
</tbody>
</table>

Table 4.10: Value of the Ly\(\alpha\) intensity calculated over a region in the flare footpoints for both flares: we calculated the values of the M1.4 flare of 08 September 1999 [Rubio da Costa et al. 2009] taking the same region size and threshold range than the M6.6 flare.

4.6.6 The jet evolution

The analysis of the H\(\alpha\) and TRACE images showed the appearance of a jet-like structure at the western boundary of the active region, at position \(\sim\) [220, 590]
4.6. FLARE OCCURRED ON 28 FEBRUARY 1999

4.6.7 Study of the Magnetic Configuration

In Fig. 4.28(a), we plotted the BBSO line-of-sight magnetogram over the WL image nearest in time: we can see that the pores in the eastern part of the active region are characterized by negative polarity and that the main sunspot has a δ configuration, with both positive and negative magnetic polarities inside the same penumbra.

Overplotting the BBSO line-of-sight magnetogram over the high resolution Hα image nearest in time, we see in Fig. 4.28(b) that filament 1, that firstly becomes...
activated, is situated over the photospheric inversion line running along the δ spot.

Plotting the ± 200, 1000 G isocontours of MDI magnetogram acquired at 16:03:02 UT before the flare, over the TRACE 171 Å image acquired at 16:30:15 UT, (taking into account the rotation of the Sun⁴), we can see in Fig. 4.29 that the main negative magnetic polarity is situated in the eastern bright region, where higher coronal loops are anchored; the main positive magnetic polarity is situated in part over the middle bright region, where the small loops are anchored and in part in the most western boundary.

In this image we can also see a bright small loop connecting a knot of negative polarity with a diffuse area of positive polarity. The negative knot moved eastward due to shearing motions during the active region evolution.

We studied the magnetic topology of the active region at the coronal level using the longitudinal component of the photospheric magnetic field provided by MDI and using an extrapolation method introduced by Alissandrakis (1981), which allows us to reconstruct the 3D magnetic field above the photosphere (boundary). This method assumes that the magnetic field is force-free not only in the corona but also in the lower levels and that it vanishes at infinity.

Using a program developed by Georgoulis & LaBonte (2007), we computed the potential field (i.e. α = 0) for the magnetograms taken at 14:27 UT (Fig. 4.30(a)) on 27 February, at 00:03 UT (Fig. 4.30(b)), 16:03 UT (Fig. 4.30(c)) and 17:36 UT (Fig. 4.30(d)) on 28 February. Even if the potential field approximation is not the most realistic one, because we are assuming that the magnetic field is constant, it allows us to have some hints on how the magnetic polarities are connected before and after the flare.

⁴There are 27 minutes of difference between the MDI and TRACE observations, but because the magnetic field does not drastically change in this short period, it is possible to assume that the temporal gap is not significant.
4.6. FLARE OCCURRED ON 28 FEBRUARY 1999

Figure 4.28: Isocontours of V Stokes component acquired at BBSO on 28 February at 17:50:31 over (a): the WL image taken at 17:49:52 UT at BBSO; (b) the Hα image taken at 17:49:48 UT at BBSO. The red contours indicate the positive magnetic field at 200, 1000 counts and the blue contours indicate the negative magnetic field at 200, 1000 counts.

From the extrapolation taken about one day before the event (Fig. 4.30(a)) we can see that there are different systems of field lines, indicating the presence of higher and lower loops. In particular, we can identify four different systems of field lines: a higher arcade (red lines) connecting the easternmost (negative) and the westernmost (positive) polarities, a lower arcade (yellow lines) connecting the more diffused magnetic concentrations of both polarities located in the centre of

Figure 4.29: Isocontours of the magnetic field strength deduced from the MDI magnetogram taken at 16:03:02 over the TRACE 171 Å image taken at 16:30:15 UT. The green contours indicate the positive magnetic field at 200, 1000 G and the blue contours indicate the negative magnetic field at -200, -1000 G. The arrow shows a small bright loop which disappears during the flare.
the active region, another bundle of field lines (magenta lines) connecting the most intense negative polarity concentration with the most intense positive one, and some small loops (blue lines) connecting two emerging negative polarities and the main positive concentration in the south-west part of the active region, where the \( \delta \) spot appears in the following hours.

Later these two negative emerging polarities moved eastward, as we can see in Fig. 4.30(b), where we can distinguish the blue field lines that connect the same magnetic concentrations, but where they appear more stretched due to shearing motions. In this representation of the extrapolation we also report a green bundle of field lines that is almost parallel to the solar surface. We note that some of these field lines have the same location and recall the same shape of the loops observed at 171 Å (see Fig. 4.15(e)).

When the \( \delta \) spot appears, we can see from the connectivity deduced in the extrapolations that in the active region there is a rearrangement of the magnetic configuration in its south-west part. In fact, few minutes before the flare (Fig. 4.30(c)) we can see that a new system of field lines corresponding to the \( \delta \) configuration appeared and that the two negative polarities are in part still connected to the main positive polarity and in part are now connected to the more diffused positive concentrations in the centre of the active region (in this representation we did not report the magenta field lines, in order to avoid some confusion). Finally, after the flare the potential field extrapolation (Fig. 4.30(d)) shows that the two negative concentrations are connected only to the positive concentrations in the core of the active region and their link to the main positive spot seems to have disappeared.
4.6. FLARE OCCURRED ON 28 FEBRUARY 1999

We studied also the variation of the positive and negative magnetic flux (using MDI/SOHO data, with the same FOV as in Fig. 4.31: 289 x 289 arcsec around the active region) from the day before the flare (27-Feb-1999) to the day after the flare (01-March-1999). We can see (Fig. 4.31) that the magnetic flux is increasing till a maximum and then it remains constant. The green vertical line indicates the time of the M6.6 flare occurrence. The red values indicate the positive magnetic flux and the blue values the negative magnetic flux.

4.6.8 Discussion on the results obtained from the study of the flare occurred on 28 February 1999

We studied the M6.6 flare using multi-wavelength observations that allowed us to follow its evolution from the photosphere to the corona. Moreover, both BBSO and MDI magnetograms were analyzed in order to provide information on the active region magnetic configuration.

The results obtained from the analysis of TRACE and BBSO data demonstrate the occurrence of a sequence of phenomena that can be described in terms of a scenario where a first filament destabilization acts as a trigger for a two-ribbon flare and in about fifteen minutes causes the activation of a second filament belonging to the same active region.

More precisely, initially a filament (filament 1) is activated in the most western side of the active region. This filament is located along the region separating the opposite polarities of a δ spot.

The bright wave-like feature observed both in the chromosphere and the corona reflects the morphology of filament 1, indicating that the plasma forming the filament is heated. This destabilization may occur as a result of internal reconnection in the filament or interaction with surrounding field. In particular, we note that a small loop joining the central negative polarity and the main positive polarity close to the filament, rises up and appears to touch the filament just at the time of brightening. The activity in this small loop indicates that the central negative polarities are also affected by the filament activation, providing a possible link to the eastern part of the active region. One of these negative polarities is located at the end of filament.
2, so reconfiguration of the linked field may result in the observed brightening and spray. The effects are dramatic because of the high shear in the core of the active region.

The associated brightness increase, having initially this wave-like shape, broadens along the magnetic inversion line and, while assuming the shape of two ribbons, also propagates toward the eastern part of the active region. Here a second filament (filament 2) is activated approximately three minutes before the flare soft X-ray peak and subsequently develops on a larger area, giving rise to a spray which involves also a quiet Sun area in the south-east part of the active region, propagating along the border of a supergranular cell. A CME was observed by LASCO/C2 coronograph at 17:54 UT and, according to the height-time measurements, the CME was initiated at around 16:42 UT, that is shortly after the activation of the second filament.

Taking account this sequence of events and the results obtained by the magnetic field extrapolation, we can infer that the second filament destabilization was a consequence of the re-arrangement of the magnetic field line connectivity caused by reconnection occurring in the magnetic field separator over the δ spot. In particular, the comparison between the magnetic field extrapolation before and after the flare shows a completely different morphology for the lower arcades (green and yellow lines in Fig. 4.30(a) and (b)) of the active region, therefore confirming the occurrence of the reconnection process. Moreover, the trend of both positive and negative magnetic flux shown in Fig. 4.31 indicates that no significant flux emergence or disappearance was observed at the time of the flare occurrence, indicating that the change in the magnetic field morphology can be ascribed to the occurrence of a reconnection process.

The analysis of TRACE data has also shown a jet-like structure at the western boundary of the active region starting before the beginning of the flare. This jet-like emission has an intermittent behavior and takes place several times during the flare. The shape of the jet-like emission resembles the structure often observed when reconnection at a null point occurs, but unfortunately the data we had did not allow us to further analyze this structure neither to single out its role in the flare event.

The TRACE data available at 1600 and 1216 Å during the peak of the flare were used to determine the power at the Lyα footpoints \( (8.1 \times 10^{25} \text{ erg s}^{-1}) \) and a comparison with the value obtained in a previous work for an M1.4 flare \( (2.4 \times 10^{25} \text{ erg s}^{-1}) \) (Rubio da Costa et al. (2009)), indicates that in the M6.6 flare the power emitted is greater than in the M1.4 flare.
Chapter 5

Study of a Solar Flare in Hα

5.1 The Hα line profile during solar flares

Hydrogen is the most abundant element in the Sun (≈ 70%). Between all the possible transitions (See Fig. 5.1), we will consider in this chapter the Balmer series and more precisely the transition at 6563 Å between the \( n = 3 \) and \( n = 2 \) energy levels, called Hα line.

When observing the Sun with ground-based instruments, a Hα filter is commonly used to study the solar chromosphere. The Hα line is a strong line, which appears in absorption in the quiet Sun or active region plages, but in emission during solar flares; e.g. Figs. 5.2 from Canfield et al. (1990).

When observing in the Hα line, it should be taken into account that the emission comes from the photosphere in the wings and from the chromosphere in the center of the line (See Fig. 1.2).

5.2 A C7.3 solar flare observed in Hα

We study a C7.3 class flare which occurred on 03 July 2002, starting at 11:43 UT in active region NOAA 10019 (S18 E21). The maximum was detected at 11:54 UT in X-ray and two minutes later, at 11:56 UT in the optical range; it lasted around 39

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1The Balmer series is due to electron transitions from higher energy levels into the \( n = 2 \) orbit; the corresponding wavelengths belong to the visible range of the electromagnetic spectrum.

Figure 5.1: The principal hydrogen spectral series. The wavelengths are given in nm units. Image taken from http://library.thinkquest.org/C006669/media/Chem/img/Series.gif
Figure 5.2: (a): $\text{H}\alpha$ line profile evolution during the 23 June 1980 flare. The pre-flare profile has been already subtracted. The Doppler shift of the line center is indicated with a bisector line. (Canfield et al. 1990); (b): Temporal variation of the $\text{H}\alpha$ intensity profile during the 20 June 1982 flare. (Ichimoto & Kurokawa 1984).

minutes in the optical range, until 12:23 UT and till 12:04 UT in X-rays (See Fig. 5.3).

Figure 5.3: GOES 10 flux as a function of time measured on 03 July 2002. The data were taken every three seconds.
The active region was observed in different wavelengths by distinct observatories before, during and after the occurrence of the flare:

- Kanzelhöle Solar Observatory, in Austria: full disk Hα images taken before the flare, at 07:07:56 UT. (See Section 5.2.2.1).
- Nancay Radiotelescopograph (NRH) at the Meudon Observatory: full disk radio data, taken at the beginning of the gradual phase of the flare, at 12:10:03 UT at 164 MHz and after the flare, at 13:40:03 UT at 327 MHz (See Table 5.1). More details in Section 5.2.2.2.

<table>
<thead>
<tr>
<th>Frequency</th>
<th>Time Range (UT)</th>
</tr>
</thead>
<tbody>
<tr>
<td>164 MHz</td>
<td>09:10:04; 10:40:03; 12:10:03; 13:40:03</td>
</tr>
<tr>
<td>327 MHz</td>
<td>09:10:04; 13:40:03</td>
</tr>
</tbody>
</table>

Table 5.1: Radio data available from Nancay Radiotelescopograph on 03 July 2002 at 164 and 327 MHz.

- SOHO/MDI: magnetograms are available before (from 01:39 to 11:15 UT, every ~ 90 minutes) and after the flare (from 12:51 to 22:23 UT, every ~ 90 minutes). As the active region was situated near the limb of the Sun, the value of the component of the magnetic field along the line of sight did not correspond to the longitudinal component of the magnetic field. The SOHO/MDI data are analyzed in Section 5.2.2.3.
- Ondřejov Observatory: MFS Hα data (see Fig. 5.4), acquired during the impulsive and gradual phase, from 11:56:24 to 12:10:42 UT. (See Section 5.2.3.1).
- TRACE: 171, 1216, 1550, 1600 Å and white light images covering the flare timing; there are also some 1700 Å data available after the flare. (See Table 5.2). (See Section 5.2.3.2).

<table>
<thead>
<tr>
<th>Filter</th>
<th>Time Range (UT)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1216</td>
<td>10:04:26 - 13:32:10</td>
</tr>
<tr>
<td>1550</td>
<td>10:04:17 - 13:32:01</td>
</tr>
<tr>
<td>1600</td>
<td>10:04:04 - 13:59:57</td>
</tr>
<tr>
<td>171</td>
<td>10:00:51 - 13:40:00</td>
</tr>
<tr>
<td>White Light</td>
<td>10:00:51 - 13:40:00</td>
</tr>
</tbody>
</table>

Table 5.2: TRACE data available for the flare occurred on 03 July 2002.

- RHESSI: X-ray data acquired between 11:50:14 and 12:05:56 UT. (See Section 5.2.3.3).
- BBSO: Hα full disk images taken after the flare occurrence, from 16:10:36 to 23:53:24 UT, with a 15 cm Singer Telescope, with spatial resolution of 1.05 arcsec/pixel. There are also CaII K and white light full disk images at 16:52:43 and 16:36:24 UT respectively. (See Section 5.2.4.1).

5.2.1 Evolution of the Active Region NOAA 10019

The active region appeared on 29 June 2002 as a beta region coming over the east limb. It was observed for 12 days and it had associated some flare events (BBSO Solar Activity Report). The evolution of the active region, deduced by BBSO Solar Activity Report, is shown in Table 5.3.
Figure 5.4: Zoom of the active region NOAA 10019. The full disk and high resolution images were taken at the Ondřejov Observatory at 12:04:21 and 12:03:12 UT respectively, during the gradual phase of the flare, indicated by an arrow.

<table>
<thead>
<tr>
<th>Day</th>
<th>Position</th>
<th>Magnetic Configuration</th>
<th>Activity</th>
</tr>
</thead>
<tbody>
<tr>
<td>29 June 2002</td>
<td>S17 E78</td>
<td>β class</td>
<td>Appearing from east limb</td>
</tr>
<tr>
<td>30 June 2002</td>
<td>S17 E62</td>
<td>βγδ class</td>
<td>C class events</td>
</tr>
<tr>
<td>01 July 2002</td>
<td>S18 E51</td>
<td>βγδ class</td>
<td>Region unchanged</td>
</tr>
<tr>
<td>02 July 2002</td>
<td>S18 E37</td>
<td>βγδ class</td>
<td>Chance for C-class events</td>
</tr>
<tr>
<td>03 July 2002</td>
<td>S18 E23</td>
<td>βγδ class</td>
<td>Chance for C-class events</td>
</tr>
<tr>
<td>04 July 2002</td>
<td>S18 E10</td>
<td>βγδ class</td>
<td>New flux emerging near south sunspots, rest decayed</td>
</tr>
<tr>
<td>05 July 2002</td>
<td>S17 W03</td>
<td>βγ class</td>
<td>South sunspots decayed, δ configuration disappeared</td>
</tr>
<tr>
<td>06 July 2002</td>
<td>S18 W18</td>
<td>βγ class</td>
<td>Unchanged in sunspot area</td>
</tr>
<tr>
<td>07 July 2002</td>
<td>S17 W31</td>
<td>βγ class</td>
<td>Magnetic complexity decayed</td>
</tr>
<tr>
<td>08 July 2002</td>
<td>S19 W41</td>
<td>βγ class</td>
<td>Slight decay</td>
</tr>
<tr>
<td>09 July 2002</td>
<td>S18 W71</td>
<td>β class</td>
<td>Decaying phase</td>
</tr>
<tr>
<td>10 July 2002</td>
<td>S19 W80</td>
<td>β class</td>
<td>Decaying phase, region at west limb</td>
</tr>
</tbody>
</table>

Table 5.3: Evolution of the active region NOAA 10019, reported by BBSO.

5.2.2 Before the flare occurrence

We studied the morphology of the active region before the flare by using the Hα images acquired at Kanzelhöhe Solar Observatory and the radio data from Nancay Radioheliograph. The magnetograms from SOHO/MDI were used to study the magnetic configuration of the active region.

5.2.2.1 Kanzelhöhe Observatory

On 03 July 2002, the telescope acquired some Hα full disk images before the flare occurrence. As I mentioned in Section 3.7, the Hα images have a spatial resolution of
5.2. A 7.3 SOLAR FLARE OBSERVED IN Hα

5.2.2 Nancay Radioheliograph

Fig. 5.6(a) shows the emission of the full disk sun at 164 MHz at 12:10 UT, during the gradual phase of the flare. As I already mentioned in Section 3.8, the spatial resolution is very low: 6 arcsec/pixel, but it is possible to recognize a brighter feature in the South-Eastern part, most probably related with the emission of the flare in NOAA 10019. Fig. 5.6(b) shows the full sun emission at 13:40 UT, when the brightening related with the flare has already disappeared.

Figure 5.5: (a): Hα image taken at the Observatory of Kanzelhöhe, showing the active region NOAA 10019 at 07:07:56 UT, before the flare; (b): MDI ± 550 Gauss isocontours at 06:24:00 UT over the Hα image taken at 07:07:56 UT.

2.23 arcsec/pixel. Fig. 5.5 shows the active region at 07:07:56 UT; the arrow shows a small dark filament between the positive and negative magnetic polarities.
5.2.2.3 SOHO/MDI magnetograms

Fig. 5.7 shows the configuration of the magnetic field in the active region two hours before the flare, at 09:39 UT (Fig. 5.7(a)) and at 11:15 UT, some minutes before the flare (Fig. 5.7(b)). Comparing Fig. 5.7(a) with Fig. 5.7(b), it is possible to infer that the positive magnetic field has decreased in the western part, in the region indicated by the arrow.

In Fig. 5.8 the ± 550 G isocontours relevant to the MDI magnetogram acquired at 11:15:00 UT, are overlayed over the TRACE and H\(\alpha\) images closer in time. Fig. 5.8(a) shows that the smaller sunspot associated to the flare is situated over a negative polarity and the most western positive polarity is situated over a small dark region. Fig. 5.8(c) shows some coronal loops connecting both positive and negative polarities. In Fig. 5.8(b) we can see that at 1216 Å the brightness increase is between both polarities, closer to the positive one. Fig. 5.8(d) shows the magnetic polarities in the H\(\alpha\) image; the arrow indicates the filament.

As already mentioned, the active region is near the solar limb and it is not possible to do an extrapolation of its magnetic field.

5.2.3 During the flare occurrence

5.2.3.1 Ondřejov Observatory

By using the H\(\alpha\) data taken with the Multichannel Flare Spectrograph (MFS), from the Ondřejov Observatory, we estimated the H\(\alpha\) intensity during the end of the impulsive phase and the gradual phase of the flare.

As explained in Section 3.5.1, the MFS takes images of the solar chromosphere at different wavelengths at the same time. On 03 July 2002, it observed the C7.3 class flare during its gradual phase in H\(\beta\), H\(\alpha\), Ca II 8542 Å and with slit-jaw. Fig. 5.9 shows an example of the data taken with the MFS at 11:57:42 UT.

Data Analysis  The times of the data taken in H\(\alpha\) during the flare are shown in Table 5.3. There are also simultaneous data relevant to a quiet region in order to
5.2. A C7.3 SOLAR FLARE OBSERVED IN Hα

Figure 5.8: MDI ±550 G isocontours taken at 11:15:00 UT over (a): TRACE WL at 11:08:56 UT; (b): TRACE 1216 Å at 11:09:02 UT; (c): TRACE 171 Å at 11:16:17 UT; (d): Hα at 11:57:39 UT. The blue contours indicate the negative magnetic field at -550 G and the red ones, the positive magnetic field at +550 G.

Table 5.4: Ondřejov Hα data relevant to active region NOAA 10019, acquired during the flare on 3 July 2002.

<table>
<thead>
<tr>
<th>Time (UT)</th>
<th>11:56:24</th>
<th>11:59:09</th>
<th>12:00:51</th>
<th>12:03:04</th>
</tr>
</thead>
<tbody>
<tr>
<td>11:56:49</td>
<td>11:59:29</td>
<td>12:01:12</td>
<td>12:03:59</td>
<td></td>
</tr>
<tr>
<td>11:58:24</td>
<td>12:00:06</td>
<td>12:02:03</td>
<td>12:10:42</td>
<td></td>
</tr>
<tr>
<td>11:58:43</td>
<td>12:00:20</td>
<td>12:02:25</td>
<td></td>
<td></td>
</tr>
</tbody>
</table>

The instrument provides fits files of calibrated Hα line profiles I(λ), where the wavelength (λ) units are Å and the intensity (I), in erg sr⁻¹ cm⁻² s⁻¹ Hz⁻¹. All profiles were obtained as an average profile from a box of 7 pixels width.

Because the active region is near the solar limb, we should take into account the
limb darkening. This effect is being corrected taking into account the method of Allen (1973), in which it was measured the refractive index of solar atmosphere as function of distance r from the solar center.

Results Having the intensity (already corrected from the quiet sun) at different wavelengths along the Hα line profile, from the center to ±2.5 Å and at different times, we studied how the line profile changes in time.

Fig. 5.10 shows the evolution of the Hα line profile, already corrected of quiet sun, during the gradual phase of the C7.3 class flare, from 11:56:24 to 12:10:42 UT. We notice that the line is in emission, as we expected. At the impulsive phase, the Hα line profile has a main emission peak; later, the profile shows secondary emission peaks at the wings of the line (see for example, Fig. 5.10(f) and (k)), most probably due to the emission of the heated background plasma at the chromosphere.

Integrating along the line profile, for a width of Δλ = ±2.5 Å from the line centre, we determined the Hα light curve, shown in Fig. 5.11. The procedure of calibration and processing of each image is divided by different steps which introduce uncertainties. It is not easy to store the parameters of all the steps. According to the experience of different observers, the final error is lower than the 15-20%.

The plot shows that the intensity decrease during the gradual phase, having many spikes from 11:56:16 to 12:01:50. At 12:01:10 there is an abrupt minimum, most probably due to the movement of the slit, which was not centered in the flare region for a few seconds. At 12:01:50 the intensity starts to decrease from 3×10^6 to 1.8×10^6 erg s^{-1}cm^{-2}sr^{-1}.

Fig. 5.12 shows a sequence of Hα images during the gradual phase, from 11:57 to 12:01 UT. The blue contours show isolevels at 25 and 90% of the maximum intensity, surrounding the eastern and bigger sunspots and the emission of the flare; the arrow shows the position of the filament, which has evolved, showing variations

\[\text{2}^\text{The limb darkening is due to the decrease of temperature of the photosphere with height: As the light we observe comes from a constant optical depth, the limb radiation is characteristically cooler than from the center of the sun and thus will be less intense (and slightly redder) than that from the center.}\]

\[\text{3}^\text{In Chapter II we will see that the Hα emission at the centre of the line comes from the photosphere and from the chromosphere at the wings.}\]
5.2. A C7.3 SOLAR FLARE OBSERVED IN Hα

Figure 5.10: Evolution of the Hα line profile, already corrected of quiet sun, during the gradual phase of the C7.3 class flare.

in brightness with respect to 07:07:56 UT (Fig. 5.5). At the beginning of the gradual phase (Fig. 5.12(a)) an increased brightness area is visible at the flare region, which later develops enclosing the filament region (See Fig. 5.12(b) and (c)). Fig. 5.12(d) shows that at 12:01 the emission starts to decrease.

5.2.3.2 TRACE data

Table 5.2 shows the TRACE data available on 03 July 2002 during the flare. The image size is 768 × 768 pixels, taken approximately every 10 minutes, except the 171 Å images, which have a time cadence of ~ 50 seconds. Some TRACE 1700 Å data are available after the flare.

In Fig. 5.13 (a), Hα isocontours are overlayed over the TRACE 1216 Å image acquired at 11:55:49 UT. Fig. 5.13(b) shows the TRACE WL image, where, besides
the main sunspots in the eastern side, we also distinguish some smaller sunspots in the western part; the H\(\alpha\) isocontours show that the flare emission is associated with these small sunspots. Observing in detail Fig. 5.13(b), there is some emission in white light associated to the flare at [-345, -355] arcsec, in the middle of the H\(\alpha\) isocontours, pointed by the arrow, which means that the flare emits also in the photosphere.

Studying the evolution of the flare at 171 Å, we observe that from 11:41:31 UT (Fig. 5.14(a)) to 11:43:07 UT (Fig. 5.14(b)), there is an increase of the brightness in the region [-325:-360,-340:-365], becoming more important in time. At 11:45:09 UT (Fig. 5.14(c)) the brightness increases through the south-eastern part of the region. According to Fig. 5.17, the flare has its maximum in 171 Å at 11:59:53 UT (Fig. 5.14(d)). At 12:01:26 UT (Fig. 5.14(e)) there are some small and compact post-flare loops, that become brighter at 12:05:31 UT.

Fig. 5.15 shows the flare at its maximum in the chromosphere and corona. In Fig. 5.15(b), the image at 1216 Å, shows that in the high chromosphere, the flaring region covers an area greater than at 171 Å.

Even if the time cadence in this wavelength is different than at 171 Å (10 minutes instead than 50 seconds), we deduced that there is an increase of brightness in the same region like in 171 Å at 11:42:03 UT (Fig. 5.16(a)), being maximum at 11:55:49 UT (Fig. 5.16(b)) and disappearing almost completely at 12:16:34 UT (Fig. 5.16(c)).

For the other wavelengths (1600 Å and 1550 Å) the evolution of the flare is quite similar to the one observed at 1216 Å.

Choosing a region of about 75 arcsec around the flare footpoints in the image already normalized in time, we calculated the total flux in DN s\(^{-1}\) at different times. Fig. 5.17 shows the variation of the flux in time for 171 Å, 1216 Å, 1600 Å and 1550 Å. The maximum of the flare happened almost 2 minutes before in 1600 Å than in 171 Å. The peak is almost simultaneous at 1216, 1600 and 1550 Å.
5.2. A C7.3 SOLAR FLARE OBSERVED IN Hα

Figure 5.12: Hα images during the gradual phase of the flare. The blue contours show isolevels at 25 and 90% of the maximum of the intensity, surrounding the western and bigger sunspots and the emission of the flare. The arrow shows the position of the filament.

5.2.3.3 RHESSI data

RHESSI entered in flare mode at 11:51:30 UT and at 12:15:40 the Sun was occulted by the Earth: we have therefore X-ray data from the end of the impulsive phase to the gradual phase of the flare.

Fig. 5.18 shows the RHESSI emission between 12 and 25 keV using two different algorithms: Pixel and Clean. In this case the most convenient is the Pixon algorithm (Fig. 5.18(b)), which managed to subtract the noise, leaving only the emission of the flare at this energy range.

Overplotting the RHESSI 12-15 keV contours over the image acquired by TRACE at 1216 Å closer in time (Fig. 5.19(a)), we see that the chromospheric brightness is well correlated with the X-ray emission at this energy range. The same behavior happens at the corona, (TRACE 171 Å; Fig. 5.19(b)). Fig. 5.19(c) shows the 12-25 keV contours over the Hα image, exhibiting a similar behavior.
5.2.4 After the flare occurrence

5.2.4.1 Big Bear Solar Observatory

The BBSO Hα images were taken with a 15 cm Singer Telescope, from 16:10 to 23:53 UT, with a time cadence of ~ 30 minutes.

In Fig. 5.20, we can see the evolution of the active region after the flare. At 16:10 UT, there is a small filament in the position [-310,-350], indicated with an arrow in
5.2. A C7.3 SOLAR FLARE OBSERVED IN Hα

Figure 5.15: TRACE images at different wavelengths at the maximum of the flare (11:55 UT).

Fig. 5.20(a) associated with the flare eruption. In the following hours the filament seems to become longer and darker (Figs. 5.20(b) and (c)).

Fig. 5.21 shows the MDI ±550 G isocontours over the BBSO Hα images closer in time, at 16:10, 19:30 and 22:59 UT respectively. The images show again that the filament is situated between the positive and negative polarities associated to the

Figure 5.16: Evolution of the flare at 1216 Å.
5.2.5 Conclusions

The active region NOAA 10019 shows a very small filament in \( \text{H}_\alpha \) (Fig. 5.5) which is activated during the flare (Fig. 5.12). After the flare eruption, there is a destabilization of the magnetic field around the filament which provokes a change in its morphology, becoming much bigger in time as is shown in Fig. 5.20.
5.2. A C7.3 SOLAR FLARE OBSERVED IN $\text{H} \alpha$

Figure 5.19: RHESSI 30 and 50% contours at 11:57:36 UT using the Pixon algorithm over (a): TRACE 1216 Å at 11:55:49 UT; (b): TRACE 171 Å at 11:57:28 UT; (c): $\text{H} \alpha$ at 11:57:39 UT.

Figure 5.20: $\text{H} \alpha$ full disk images taken with the 15 cm Big Bear telescope, zoomed at the active region NOAA 10019 at different times after the flare.

Comparing the variation of the $\text{H} \alpha$ intensity (Fig. 5.11) with the X-ray flux from GOES 0.5-4 Å (Fig. 5.3), we see that the flux in X-ray starts to decay earlier than in $\text{H} \alpha$ and with a higher slope.

In the next Chapters 6 and 7, I will study the $\text{H} \alpha$ emission during solar flares.

Figure 5.21: MDI $\pm$550 G isocontours over the closest $\text{H} \alpha$ image in time. (a): MDI contours taken at 16:03:00 UT over the $\text{H} \alpha$ image taken at 16:10:36 UT; (b): MDI contours taken at 19:30: UT over the $\text{H} \alpha$ image taken at 19:30:11 UT; (c): MDI contours taken at 22:23:00 UT over the $\text{H} \alpha$ image taken at 22:59:24 UT.
theoretically in order to compare the intensity, the line profile and the time evolution obtained by simulations with the observations shown in this Chapter.
SIMULATIONS
Chapter 6

Study of lines spectrum of the Hydrogen atom

6.1 Introduction

We used a radiative transfer code developed initially by Gouttebroze et al. (1978) in order to compute the spectral lines profiles emitted by the hydrogen atom and compare them with the observations.

In Section 5 we presented the study of a solar flare in Hα for which there are spectroscopic data. With the radiative transfer code, we can compare the simulated Hα intensity with the observed one.

For the flares in Ly-α, we do not have the detailed line profile. In order to compare the intensity we estimated at the flare footpoints (see Section 4.6.5.1) with the results from the simulations, we integrated the emergent intensity along the computed Lyα line profile (see Section 6.4.1.1). In this chapter we computed the line profiles corresponding to various solar atmospheric models from Machado et al. (1980): the quiet sun, the plage and bright and faint flares. Changing the temperature and microturbulent velocity profile along the solar atmosphere, we studied how these parameters affect the hydrogen line profiles and their total integrated intensity.

6.2 The code

The basic method of solution is described in detail in Gouttebroze et al. (1978); in this Section we will briefly outline the procedure: assuming a twenty-level and continuum hydrogen model atom, the code solves the equations of population conservation and radiative transfer for a one-dimensional plane-parallel atmosphere out of Local Thermodynamic Equilibrium (LTE). It produces, as output parameter, the intensity as a function of frequency and depth for the lines and continua emitted by the hydrogen atom, including the Lyman and Balmer se-
The program gives also the possibility to study the variation of the intensity with the angle-of-sight, from the centre of the sun to the solar limb, changing the value of $\mu = \cos \theta$, where $\theta$ is the angle between the radial direction and the line of sight, as Fig. 6.1 displays.

### 6.2.1 The Radiative Transfer Equation

The general transport equation is given by Eq. 6.1 (Mihalas 1978; Shu 1991):

$$\frac{dI_\nu}{ds} = j_\nu - \alpha_\nu I_\nu,$$

(6.1)

where $j_\nu$ is the monochromatic emissivity and $\alpha_\nu$, the extinction coefficient.

Assuming a plane-parallel atmosphere\(^1\), and taking into account that the source function, $S_\nu = \frac{j_\nu}{\alpha_\nu}$, is the emissivity divided by the extinction coefficient and that the optical depth, $\tau_\nu = -\alpha_\nu ds$, can be defined as the density times the monochromatic mass extinction coefficient, the Eq. 6.1 can be written as Eq. 6.2:

$$\mu \frac{dI_\nu}{d\tau_\nu} = I_\nu - S_\nu$$

(6.2)

Applying as boundary conditions that there is no irradiation from outside: $I_\nu(0, \mu) = 0$ and that for $\tau_\nu \to \infty$, $S_\nu(\tau_\nu) e^{-\tau_\nu} \to 0$, the solution of the transfer equation is given by Eq. 6.5 (Chandrasekhar 1960; Gray 1992):

$$I_\nu(\tau_\nu, +\mu) = \int_{\tau_\nu}^{\infty} S_\nu(t_\nu) e^{-(t_\nu-\tau_\nu)/\mu} dt_\nu / \mu$$

(6.3)

$$I_\nu(\tau_\nu, -\mu) = \int_{0}^{\tau_\nu} S_\nu(t_\nu) e^{-(t_\nu-\tau_\nu)/\mu} dt_\nu / |\mu|$$

(6.4)

Gouttebroze et al. (1978) solve the radiative transfer equation using the Feautrier’s method (Feautrier 1964) with variable Eddington factors (Auer & Mihalas 1970).

### 6.2.2 The Statistical Equilibrium Equation

The statistical equilibrium means that the level populations do not vary in time. This is expressed in Eq. 6.6

$$\frac{dn_i(\vec{r})}{dt} = \sum_{j \neq i}^{N} n_j(\vec{r}) P_{ji}(\vec{r}) - n_i(\vec{r}) \sum_{j \neq i}^{N} P_{ij}(\vec{r}) = 0,$$

(6.6)

where $n_i$ is the population of the level $i$, $N$ is the total number of levels and $P_{ij}$ is the transition rate between the levels $i$ and $j$, which is defined as the sum of the radiative processes plus the collisional processes ($P_{ij} = R_{ij} + C_{ij}$).

\(^1\)It assumes that the intensity depends only on one space coordinate, the vertical depth.
6.3. THE ATMOSPHERIC MODELS

6.2.3 The Hydrostatic Equilibrium Equation

Assuming a stationary atmosphere, we get that Eq. 6.7 is satisfied:

\[ \frac{dP}{dz} = -g \rho, \]  

(6.7)

where \( z \) is the geometrical depth.

Assuming that the contributions to the pressure, \( P \) is the sum of the atomic, electronic and microturbulent pressures, it can be written as in Eq. 6.8:

\[ P = -\left( \frac{c_2 + \eta}{c_1} \cdot \kappa T + 0.5 \cdot v_T^2 \right) \frac{dn}{dz}, \]  

(6.8)

where \( c_1 = \sum_i A_i m_i \), being \( A_i \) the abundance of the element \( i \) with respect to \( H \) and \( m_i \) the atomic mass of \( i \); \( c_2 = \sum_i A_i \) and \( \eta = N_e/N_H \), with \( N_H \) the hydrogen population and \( N_e \) the electron population and \( v_T \) is the microturbulent velocity.

Substituting Eq. 6.8 into Eq. 6.7 we get Eq. 6.9 to calculate the altitudes of the grid points:

\[ z - z_0 = \int_{n_H(z_0)}^{n_H} \left( \frac{c_2 + \eta}{c_1} \kappa T + 0.5 \cdot c_1 \cdot v_T^2 \right) \frac{dn}{dx} dx, \]  

(6.9)

where \( x \) is an integration variable (\( dx \equiv dn/H \)).

6.2.4 The method

Giving some initial values of \( \eta \) and \( z \) and using the Eq. 6.9 iteratively, the code gets new values of \( z \), \( N_e \) and \( N_H \), which are needed to compute the line profiles and are used as input parameters for the statistical equilibrium equation and the radiative transfer equation.

Solving simultaneously the statistical equilibrium equation and the radiative transfer equation, by using the Feautrier’s method, the code obtains the intensity as a function of frequency and depth (Gouttebroze et al. 1978).

6.3 The atmospheric models

The code uses as input parameter an atmospheric model prescribing the variation of temperature and microturbulent velocity with column mass. Initially we used two temperature-height distribution and microturbulent velocity flare models given by Machado et al. (1980): F1, for faint flares and F2, for brighter flares. These are also compared with a quiet Sun model (model C of Vernazza et al. (1981)).

Fig. 6.2 shows the different temperature models. In Fig. 6.2(a) there are the four different temperature models of Basri et al. (1979): the quiet sun (model QS), plage model (model P), faint flares model (model F1) and bright flares (model F2). Fig. 6.2(b) shows in detail the bright flares (blue line) and faint flares (red line) models. They are differentiated mainly by the height of the transition region, in which the temperature changes abruptly: the quiet sun model is the model with the highest altitude transition region, which is followed by the faint flares model; bright flares temperature model is the one with a deeper transition region.

Fig. 6.3 shows the variation of the microturbulent velocity for three atmospheric models: quiet sun (QS), bright flares (F2) and faint flares (F1). For the quiet sun model (green line) the microturbulent velocity is decreasing at the corona and it
starts to increase at the chromosphere; for faint flares (red line), it starts decreasing abruptly from the corona, then it remains constant and at the chromosphere it starts to decrease again; the bright flares model (blue line) has the same behavior than the faint flares model, but it remains constant for longer and the final decay is faster.

Some different models between F1 and F2 have been obtained to study how the intensity of different lines might change with the variation of the temperature and the microturbulent velocity at several layers.

Figure 6.3: Variation of the microturbulent velocity with column mass for different models: faint flares (red line-model F1), bright flares (blue line-model F2) and quiet sun (green line-model QS).
6.3. THE ATMOSPHERIC MODELS

6.3.1 Changing the temperature

We calculated different models as linear combination between the F1 and F2 temperature models, in order to obtain different heights for the transition region.

Both models \( T_{F1} \) and \( T_{F2} \) have been interpolated using a cubic spline interpolation, obtaining the values of the temperature at each value of a generic array of column mass. Fig. 6.4 shows the available values of the models as red points and the quality of the interpolation as a green line.

6.3.1.1 Model temperature \( T_1 \)

The model \( T_1 \) is obtained as a linear combination between the models F1 and F2 from Machado et al. (1980); it was obtained calculating the average of the temperature (\( T \)) between F1 and F2, for each column mass average value (\( m \)), as it is explained in Eq. 6.10

\[
mT_1 = \frac{m_{F1} + m_{F2}}{2} \tag{6.10}
\]

\[
T_{T1} = \frac{T_{F1} + T_{F2}}{2}
\]

Fig. 6.5(a) shows the model \( T_1 \), which has a transition region situated between the one of the models F1 and F2.

Figure 6.4: The values of the temperature of the models F1 and F2 have been interpolated. The result is represented with the green line, while the red points are the values of the models from Machado et al. (1980). The figures at the bottom show a zoom to visualize the details of the fit in the transition region.
CHAPTER 6. LINES SPECTRUM OF HYDROGEN ATOM

6.3.1.2 Model temperature \( T_2 \)

In order to obtain a model with the base of the transition region located deeper in the atmosphere than \( F_2 \), we calculated model \( T_2 \), which has the same temperature values as model \( T_1 \), but for larger values of the column mass deeper transition region.

It was obtained in the following way: from the corona to the transition region, the value of the column mass \( m_{T2} \) is the sum of \( m_{F1} + m_{F2} \); after the sudden change of the temperature gradient at the base of the transition region, the resulting value of \( m_{T2} \) is the same than \( m_{T1} \). Eq. 6.11 explains mathematically how the model \( T_2 \) is obtained.

\[
m_{T2} = \frac{m_{F1} + m_{F2}}{2}, \quad \text{above the base of the transition region} \tag{6.11}
\]

\[
m_{T2} = m_{F1} + m_{F2}, \quad \text{below the base of the transition region}
\]

\[
T_{T2} = \frac{T_{F1} + T_{F2}}{2}
\]

The model \( T_2 \) is represented in Fig. 6.5(b), confirming that the base of the transition region and the base of the corona are located deeper in the atmosphere of the Sun, i.e., at a larger column mass, than the other models. The variation of the temperature in the chromosphere is between the models \( T_{T1} \) and \( T_{T2} \), allowing us to study how it affects the intensity of the hydrogen lines.
6.3. THE ATMOSPHERIC MODELS

6.3.1.3 Model temperature $T_3$

Model $T_3$ has a corona situated at a higher layer than model $F_1$. It was calculated using the same temperature values as model $T_1$, but recalculating the values of the column mass: from the corona to the transition region, the resulting column mass is 0.4 times smaller than the one used at the model $T_1$; from the transition region to deeper layers, the values of the column mass are the same than the model $T_1^2$. Eq. 6.12 explains mathematically how the model $T_3$ was obtained.

$$m_{T_3} = 0.4 \times \frac{m_{F1} + m_{F2}}{2},$$  \hspace{1cm} \text{above the base of the transition region} \hspace{1cm} (6.12)

$$m_{T_3} = \frac{m_{F1} + m_{F2}}{2},$$  \hspace{1cm} \text{below the base of the transition region}

$$T_{T_3} = \frac{T_{F1} + T_{F2}}{2},$$

Fig. 6.5(c) shows that model $T_3$ is the one having the lowest column mass in the corona and transition region, following the same variation of temperature in the chromosphere than models $T_1$ and $T_2$.

6.3.1.4 Model temperature $T_4$

In order to know how the variation of the temperature at the chromosphere affects the hydrogen lines, we created model $T_4$, which has the same temperature variation as model $T_1$ at the corona, but the temperature at the base of the transition region is lower than for model $T_1$; the variation of the temperature in the chromosphere is 0.9 times smaller than the one used in model $T_1$, as explained in Eq. 6.13.

$$m_{T_4} = \frac{m_{F1} + m_{F2}}{2}$$  \hspace{1cm} (6.13)

$$T_{T_4} = 0.9 \times \frac{T_{F1} + T_{F2}}{2},$$  \hspace{1cm} \text{above the base of the transition region}

$$T_{T_4} = \frac{T_{F1} + T_{F2}}{2},$$  \hspace{1cm} \text{below the base of the transition region}

Model $T_4$ is the one with the lowest temperature, as Fig. 6.5(d) shows.

6.3.2 Changing the microturbulent velocity

The microturbulent velocity is defined as a convolution of a Gaussian distribution with the line extinction coefficient and it is added as an extra term into the Doppler width equation (see Eq. 6.14), broadening the width of the lines profile (Rutten 2003).

$$\Delta \nu_D = \frac{\nu_0}{c} \sqrt{\frac{2\kappa T}{m} + \nu_{\text{micro}}^2}$$  \hspace{1cm} (6.14)

As in the previous section, the values of the microturbulent velocity at each value of a generic array of column mass have been obtained by interpolating the

\footnote{It was needed to extrapolate the column mass for values smaller than the model $T_{F1}$, assuming that the temperature has the same behavior till the base of the transition region}
microturbulent velocity from F1 and F2 with a cubic spline function. Fig. 6.6 shows the result.

To study how the variation of the microturbulent velocity affects the intensity and the profile of different lines, we created three different models of the microturbulent velocity, as a combination of the microturbulent velocity of faint flares ($v_{F1}$) and bright flares ($v_{F2}$).

### 6.3.2.1 Model microturbulent velocity $v_1$

The model $v_1$ has been calculated as the average between the microturbulent velocity of the models $v_{F1}$ and $v_{F2}$ for each value of column mass (See Eq. 6.15). Fig. 6.7(a) shows the resulting values.

$$v_{v1} = \frac{v_{F1} + v_{F2}}{2}$$  \hspace{1cm} (6.15)

### 6.3.2.2 Model microturbulent velocity $v_2$

As Eq. 6.16 shows, the model $v_2$ is the sum of both models $v_{F1}$ and $v_{F2}$, which is shown with a green line in Fig. 6.7(b).

$$v_{v2} = v_{F1} + v_{F2}$$  \hspace{1cm} (6.16)

### 6.3.2.3 Model microturbulent velocity $v_3$

The model $v_3$ is the difference in the microturbulent velocity value for the model $v_{F2}$ and the value of the model $v_{F1}$ at each value of the column mass, as it is explained in Eq. 6.17. (See Fig. 6.7).

$$v_{v3} = v_{F1} - v_{F2}$$  \hspace{1cm} (6.17)

The resulting values for $v_3$ and $v_2$ are unrealistic, but the aim of the study is just to observe the influence of this parameter on the intensities.

### 6.4 Results

With the radiative transfer code we studied how the profile of different hydrogen lines is affected by the temperature at different layers, according to the different values of the microturbulent velocity.

![Figure 6.6](image)

Figure 6.6: The values of the microturbulent velocity of the models F1 (Faint flares), F2 (Bright flares) and QS (Quiet Sun) have been interpolated. The fitted result is represented with the green line, while the red points are the values of the models from Machado et al. [1980].
6.4. RESULTS

Figure 6.7: The three models (green line) have been obtained like a combination between the $v_{F1}$ (red line) and $v_{F2}$ (blue line) models from Machado et al. (1980).

models shown in Fig. 6.3 and by the different microturbulent velocity models (Fig. 6.7).

At the same time, we also studied how the total integrated intensity along the line profiles changes from the centre of the Sun to the limb, with the line-of-sight angle, $\mu$.

6.4.1 Integrated intensity

We calculated the value of the integrated intensity of the Ly-$\alpha$ Ly-$\beta$ and H$\alpha$ lines profiles, for the different:

- Temperature models
- Microturbulent velocity models
- Values of the line of sight

6.4.1.1 The Ly$\alpha$ intensity

In order to know how the width of the line profile influences the total integrated intensity, we calculated it for a width of $\Delta \lambda = 2$ and 1 Å from the centre of the line. The results shown in Tables 6.1 and 6.2 demonstrate that the integrated intensity does not significantly change cutting the wings of the line at 1 Å or 2 Å, so we decided to use a width of $\Delta \lambda = 2$ Å.

<table>
<thead>
<tr>
<th>Model $T_F$</th>
<th>$\mu = 0.2$</th>
<th>$\mu = 0.6$</th>
<th>$\mu = 1.0$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$T_{F1}$</td>
<td>$2.60 \times 10^9$</td>
<td>$2.69 \times 10^9$</td>
<td>$2.77 \times 10^9$</td>
</tr>
<tr>
<td>$T_{F2}$</td>
<td>$5.40 \times 10^8$</td>
<td>$5.64 \times 10^8$</td>
<td>$5.88 \times 10^8$</td>
</tr>
<tr>
<td>$T_1$</td>
<td>$5.25 \times 10^9$</td>
<td>$5.52 \times 10^9$</td>
<td>$5.71 \times 10^9$</td>
</tr>
<tr>
<td>$T_2$</td>
<td>$7.49 \times 10^9$</td>
<td>$7.87 \times 10^9$</td>
<td>$8.18 \times 10^9$</td>
</tr>
<tr>
<td>$T_3$</td>
<td>$3.23 \times 10^9$</td>
<td>$3.38 \times 10^9$</td>
<td>$3.46 \times 10^9$</td>
</tr>
<tr>
<td>$T_4$</td>
<td>$5.54 \times 10^6$</td>
<td>$5.72 \times 10^6$</td>
<td>$5.89 \times 10^6$</td>
</tr>
</tbody>
</table>

Table 6.1: Integrated intensity in erg s$^{-1}$cm$^{-2}$sr$^{-1}$ for the Ly$\alpha$ line for a range of $\Delta \lambda = 1$ Å around the centre.
**Temperature models**  From Table 6.1, we conclude that the $\text{Ly}\alpha$ intensity decreases from the centre of the Sun to the limb for all the flare models. The strongest intensity is the one associated to the model $T_2$, followed by the model of bright flares, $T_{F2}$, from [Machado et al. (1980)]; the weakest $\text{Ly}\alpha$ intensity is the one associated with the bright flares model from [Machado et al. (1980)].

<table>
<thead>
<tr>
<th>Model</th>
<th>Intensity $\mu = 0.2$</th>
<th>Intensity $\mu = 0.6$</th>
<th>Intensity $\mu = 1.0$</th>
</tr>
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<tr>
<td>Model $T_{F1}$</td>
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</tr>
<tr>
<td>Model $T_{F2}$</td>
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<tr>
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<td>$6.10 \times 10^6$</td>
<td>$6.27 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_2$</td>
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<td>$8.53 \times 10^6$</td>
<td>$8.80 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_3$</td>
<td>$3.76 \times 10^6$</td>
<td>$3.90 \times 10^6$</td>
<td>$3.98 \times 10^6$</td>
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<tr>
<td>Model $T_4$</td>
<td>$5.89 \times 10^6$</td>
<td>$5.94 \times 10^6$</td>
<td>$6.09 \times 10^6$</td>
</tr>
</tbody>
</table>

Table 6.2: Integrated intensity in erg s$^{-1}$cm$^{-2}$sr$^{-1}$ for the $\text{Ly}\alpha$ line for a range of $\Delta \lambda = 2$ Å around the centre.

From this we conclude that the depth of the transition region is directly related with the increase of the intensity of the line: the deepest models ($T_2$ and $T_{F2}$) are the brightest in $\text{Ly-}\alpha$, followed by the $T_1$ and $T_4$ models respectively, which are the consecutive ones in depth. Model $T_4$ has a different variation of temperature in the chromosphere from model $T_1$, with $T_4$ having a lower temperature gradient than $T_1$: this difference is shown in Table 6.1 as a small decrease of the intensity in $\text{Ly}\alpha$. Model $T_3$ has a transition region situated at a higher layer than the other models and a hotter, more extended chromospheric part than model $T_{F1}$; this is the reason why the $\text{Ly}\alpha$ intensity is larger for $T_3$ than for $T_{F1}$.

**Microturbulent velocity models**  Table 6.3 shows the variation of the integrated intensity along the $\text{Ly}\alpha$ line profile for a width of $\Delta \lambda = 2$ Å around the centre for the three different microturbulent velocity models.

<table>
<thead>
<tr>
<th>Model $T$</th>
<th>Model $v$</th>
<th>Intensity $\mu = 0.2$</th>
<th>Intensity $\mu = 0.6$</th>
<th>Intensity $\mu = 1.0$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Model $T_1$</td>
<td>Model $v_1$</td>
<td>$5.90 \times 10^6$</td>
<td>$6.09 \times 10^6$</td>
<td>$6.27 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_1$</td>
<td>Model $v_2$</td>
<td>$5.35 \times 10^6$</td>
<td>$5.67 \times 10^6$</td>
<td>$5.86 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_1$</td>
<td>Model $v_3$</td>
<td>$6.10 \times 10^6$</td>
<td>$6.17 \times 10^6$</td>
<td>$6.30 \times 10^6$</td>
</tr>
<tr>
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<td>Model $v_1$</td>
<td>$8.31 \times 10^6$</td>
<td>$8.52 \times 10^6$</td>
<td>$8.78 \times 10^6$</td>
</tr>
<tr>
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<td>Model $v_2$</td>
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<td>$8.38 \times 10^6$</td>
</tr>
<tr>
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<td>Model $v_3$</td>
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<td>$8.67 \times 10^6$</td>
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</tr>
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<td>Model $v_1$</td>
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<td>$3.98 \times 10^6$</td>
</tr>
<tr>
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<td>Model $v_2$</td>
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<td>$3.64 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_3$</td>
<td>Model $v_3$</td>
<td>$3.85 \times 10^6$</td>
<td>$3.95 \times 10^6$</td>
<td>$4.00 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_4$</td>
<td>Model $v_1$</td>
<td>$5.89 \times 10^6$</td>
<td>$5.94 \times 10^6$</td>
<td>$6.08 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_4$</td>
<td>Model $v_2$</td>
<td>$5.24 \times 10^6$</td>
<td>$5.47 \times 10^6$</td>
<td>$5.63 \times 10^6$</td>
</tr>
<tr>
<td>Model $T_4$</td>
<td>Model $v_3$</td>
<td>$6.17 \times 10^6$</td>
<td>$6.06 \times 10^6$</td>
<td>$6.15 \times 10^6$</td>
</tr>
</tbody>
</table>

Table 6.3: Integrated intensity in erg s$^{-1}$cm$^{-2}$sr$^{-1}$ for the $\text{Ly}\alpha$ line for a range of $\Delta \lambda = 2$ Å around line centre.
The results do not change significantly from one model to the other because the Lyα line is optically thick and the microturbulent velocity does not affect much the resulting line profile.

6.4.1.2 The Ly-β intensity

For the Lyβ line emission, we calculated the integrated intensity for a width interval of 1 Å around the centre of the line.

**Temperature models**  Table [6.4] shows the variation of the Ly-β intensity with the temperature models.

The intensity increases from the centre of the Sun to the limb; we could expect this behaviour because the variation of the line-of-sight is, in principle, intercepting more material when we look at the limb than at the centre of the disc, but in the real situation the intensity depends also on the structure of the atmosphere, the height formation of the lines, its optical thickness, etc. In fact, the Lyα line has a different behavior than the Ly-β line.

<table>
<thead>
<tr>
<th>Model</th>
<th>Intensity</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>μ = 0.2</td>
</tr>
<tr>
<td>Model T_F1</td>
<td>1.34 × 10^5</td>
</tr>
<tr>
<td>Model T_F2</td>
<td>3.49 × 10^5</td>
</tr>
<tr>
<td>Model T_1</td>
<td>3.04 × 10^5</td>
</tr>
<tr>
<td>Model T_2</td>
<td>4.64 × 10^5</td>
</tr>
<tr>
<td>Model T_3</td>
<td>1.79 × 10^5</td>
</tr>
<tr>
<td>Model T_4</td>
<td>2.78 × 10^5</td>
</tr>
</tbody>
</table>

Table 6.4: Integrated intensity in erg s^{-1}cm^{-2}sr^{-1} for the Lyβ line for a range of Δλ = 1 Å around the centre.

**Microturbulent velocity models**  Table [6.5] shows the variation of the Ly-β intensity for various microturbulent velocity models.

As in Table [6.4], the Ly-β intensity increases from the centre of the Sun to the limb. As for the Lyα intensity, the highest intensity is the one associated to model v_3, which is v_F1 − v_F2, followed by model v_1 and the lowest Ly-β intensity is the one associated with model v_2, concluding that the intensity decreases with the increase of the microturbulent velocity. This is contrary to the first expected intuitive idea because the microturbulent velocity broadens the lines, yielding larger integrated intensities, most probably due to the contribution of other factors, like the temperature structure of the atmosphere, the height formation of the line or its optical thickness.

6.4.1.3 The Hα intensity

For the Hα line we calculated the integrated for a wavelength interval of 5 Å around the centre of the line, see Table [6.6].

**Temperature models**  As for the Lyα line, the Hα intensity decreases from the centre of the Sun to the limb as expected: the Hα line is optically thick line, and the fact that at the limb of the Sun the atmosphere intercepts more material than
at the centre of the disc should affect more optically thin lines than optically thick lines.

Table 6.5: Integrated intensity in erg s$^{-1}$cm$^{-2}$sr$^{-1}$ for the Ly-$\beta$ line for a range of $\Delta \lambda=1$ Å around the centre.

<table>
<thead>
<tr>
<th>Model $T$</th>
<th>Model $v$</th>
<th>$\mu = 0.2$</th>
<th>$\mu = 0.6$</th>
<th>$\mu = 1.0$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Model $T_1$</td>
<td>Model $v_1$</td>
<td>$2.84 \times 10^5$</td>
<td>$2.32 \times 10^5$</td>
<td>$2.16 \times 10^5$</td>
</tr>
<tr>
<td>Model $v_2$</td>
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<td>$1.76 \times 10^5$</td>
<td>$1.62 \times 10^5$</td>
<td></td>
</tr>
<tr>
<td>Model $v_3$</td>
<td>$3.19 \times 10^5$</td>
<td>$2.61 \times 10^5$</td>
<td>$2.42 \times 10^5$</td>
<td></td>
</tr>
<tr>
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<td>Model $v_1$</td>
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<td>$3.58 \times 10^5$</td>
<td>$3.33 \times 10^5$</td>
</tr>
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<td>Model $v_2$</td>
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</tr>
<tr>
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<td>$3.97 \times 10^5$</td>
<td>$3.68 \times 10^5$</td>
<td></td>
</tr>
<tr>
<td>Model $T_3$</td>
<td>Model $v_1$</td>
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<td>$1.36 \times 10^5$</td>
<td>$1.27 \times 10^5$</td>
</tr>
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<td>$1.02 \times 10^5$</td>
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</tr>
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<td>$1.56 \times 10^5$</td>
<td>$1.45 \times 10^5$</td>
<td></td>
</tr>
<tr>
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<td>Model $v_1$</td>
<td>$2.60 \times 10^5$</td>
<td>$2.04 \times 10^5$</td>
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<td>$1.52 \times 10^5$</td>
<td>$1.37 \times 10^5$</td>
<td></td>
</tr>
<tr>
<td>Model $v_3$</td>
<td>$2.94 \times 10^5$</td>
<td>$2.29 \times 10^5$</td>
<td>$2.09 \times 10^5$</td>
<td></td>
</tr>
</tbody>
</table>

Table 6.6: Integrated intensity in erg s$^{-1}$cm$^{-2}$sr$^{-1}$ for the H$\alpha$ line for a range of $\Delta \lambda=5$ Å around the line centre.

Table 6.6 shows that the highest intensity is the one associated to model $T_{F_2}$, followed by the model $T_2$, which has a deeper transition region but with slower temperature at the chromosphere. The total integrated intensity of the models $T_1$ and $T_3$ (which have the same temperature variation in the chromosphere, but different depth for the transition region) are very close to the one of $T_2$. The lowest H$\alpha$ intensity is associated with models $T_{F_1}$ and $T_4$ respectively, which are the ones with the lowest chromospheric temperature. From this we conclude that the H$\alpha$ intensity is affected mainly by the temperature model in the chromosphere, as was expected because the wings of the H$\alpha$ line are formed in this atmospheric layer (see Fig. 1.2 from Vernazza et al. (1981)).

**Microturbulent velocity models** As Table 6.7 shows, the variation of the microturbulent velocity in height does not affect the H$\alpha$ intensity, the values are very close, being slightly higher for the model $v_3$. 
6.4. RESULTS

<table>
<thead>
<tr>
<th>Model $T$</th>
<th>Model $v$</th>
<th>Intensity $\mu = 0.2$</th>
<th>Intensity $\mu = 0.6$</th>
<th>Intensity $\mu = 1.0$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Model $T_1$</td>
<td>Model $v_1$</td>
<td>$2.01 \times 10^7$</td>
<td>$2.09 \times 10^7$</td>
<td>$2.23 \times 10^7$</td>
</tr>
<tr>
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<td>$1.95 \times 10^7$</td>
<td>$2.13 \times 10^7$</td>
</tr>
<tr>
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<td>$2.23 \times 10^7$</td>
<td>$2.34 \times 10^7$</td>
</tr>
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<td>Model $v_1$</td>
<td>$2.03 \times 10^7$</td>
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<td>$2.25 \times 10^7$</td>
</tr>
<tr>
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</tr>
<tr>
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<td>$2.27 \times 10^7$</td>
<td>$2.24 \times 10^7$</td>
<td>$2.34 \times 10^7$</td>
</tr>
<tr>
<td>Model $T_3$</td>
<td>Model $v_1$</td>
<td>$2.01 \times 10^7$</td>
<td>$2.08 \times 10^7$</td>
<td>$2.22 \times 10^7$</td>
</tr>
<tr>
<td></td>
<td>Model $v_2$</td>
<td>$1.71 \times 10^7$</td>
<td>$1.93 \times 10^7$</td>
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</tr>
<tr>
<td></td>
<td>Model $v_3$</td>
<td>$2.24 \times 10^7$</td>
<td>$2.22 \times 10^7$</td>
<td>$2.33 \times 10^7$</td>
</tr>
<tr>
<td>Model $T_4$</td>
<td>Model $v_1$</td>
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<td>$1.38 \times 10^7$</td>
<td>$1.53 \times 10^7$</td>
</tr>
<tr>
<td></td>
<td>Model $v_2$</td>
<td>$1.08 \times 10^7$</td>
<td>$1.33 \times 10^7$</td>
<td>$1.49 \times 10^7$</td>
</tr>
<tr>
<td></td>
<td>Model $v_3$</td>
<td>$1.19 \times 10^7$</td>
<td>$1.42 \times 10^7$</td>
<td>$1.56 \times 10^7$</td>
</tr>
</tbody>
</table>

Table 6.7: Integrated intensity in erg s$^{-1}$cm$^{-2}$sr$^{-1}$ for the H$\alpha$ line for a range of $\Delta \lambda = 5$ Å around the line centre.

6.4.2 Lines profiles

Studying also the profile of the lines for the different temperature and microturbulent velocity models, we can get more information about the line formation mechanisms at different frequencies.

If the line core is formed at a different height in the atmosphere than the wings of the line, the different atmospheric models would change the shape of the lines profiles, allowing us to study how the models act on the hydrogen lines.

6.4.2.1 The Ly$\alpha$ line profile

The Ly$\alpha$ line is an optically thick line: the core of the line is formed in the transition region; however the wings are formed in the chromosphere (see Vernazza et al. (1981)).

In Fig. 6.8 is represented the Ly$\alpha$ line profile for the temperature model $T_2$ at different line-of-sights, from the centre of the Sun ($\mu = 1.0$) to the limb ($\mu = 0.2$), showing that the strongest one is at the Sun centre, as already mentioned.

In Fig. 6.9(a) we can compare models $T_1$ and $T_4$ (two models with the same temperature in the corona and the transition region, but a different one in the chromosphere). The temperature of model $T_4$ is lower in the chromosphere up to the base of the transition region than that of model $T_1$, resulting in the Ly$\alpha$ intensity being stronger than for $T_1$.

Fig. 6.9(b) is comparing the models $T_2$ and $T_3$, which allows us to study the influence of the position of the transition region on the Ly$\alpha$ line profile. In $T_2$ the transition region is much lower in the atmosphere than in $T_3$. As already commented, the Ly$\alpha$ intensity is stronger for $T_2$ than for $T_3$.

Fig. 6.10 shows how the microturbulent velocity models affect the Ly$\alpha$ line profile, corroborating that the intensity associated to the model $v_3$ is the highest one. The model $v_2$ is the one with the lowest Ly$\alpha$ intensity, but at the same time the profile becomes broader; this is consistent with the fact that large microturbulent
Figure 6.8: Lyα line profile obtained for the temperature model $T_3$ at different line of sights, from the centre of the Sun ($\mu = 1.0$) to the limb ($\mu = 0.2$).

Figure 6.9: Lyα line profile obtained using different temperature models at $\mu = 0.2$.

velocities add to line broadening. At the same time we can see that the intensity in the core of the Lyα line is significantly affected by the microturbulent velocity model, being lower for model $v_2$ than for $v_3$. Despite the line broadening, the integrated intensity is lower for $v_2$ than for $v_3$ (see Table 6.3).

6.4.2.2 The Hα line profile

The Hα line is an absorption line in the quiet Sun, contrary to the Lyα line. According to Fig. 6.12, it is also an optically thick line and studying how the temperature, line-of-sight and microturbulent velocity affect the line profile, we can compare the behavior of both lines.

The core of the Hα line is formed down in the chromosphere and the wings in the photosphere.

The variation of the Hα line profile at different line-of-sights is shown in Fig.
6.4. RESULTS

Figure 6.10: Lyα line profile obtained for the temperature model $T_3$ at the centre of the solar disk ($\mu = 1.0$), using different microturbulent velocities.

In this case the centre of the line and also the wings vary with the line-of-sight, being broader at the limb and narrower at the disk centre.

Fig. 6.12(a) shows the models $T_1$ and $T_4$. The temperature at the corona for both models follows the same profile and the position of the transition region is the same for both models. The model $T_4$ has a lower transition region temperature, making the main difference between $T_1$ and $T_4$, resulting in the Hα intensity being larger for $T_1$ than for $T_4$.

Figure 6.11: Hα line profile obtained for the temperature model $T_{F2}$ at different line-of-sights, from the centre of the Sun ($\mu = 1.0$) to the limb ($\mu = 0.2$).
CHAPTER 6. LINES SPECTRUM OF HYDROGEN ATOM

Figure 6.12: Hα line profile obtained using different temperature models at $\mu = 0.2$.

(a) Models $T_1$ and $T_4$

(b) Models $T_2$ and $T_3$

Figure 6.13: Hα line profile obtained for the temperature model $T_3$ at the centre of the Sun ($\mu = 1.0$), using different microturbulent velocities.

In Fig. 6.12(b) we compare the models $T_2$ and $T_3$. The transition region of the model $T_2$ is deeper than for the model $T_3$, forming a line with a more intense core; both models have the same chromospheric temperature behavior, making the wings of the Hα line almost identical for these two models.

The Fig. 6.13 shows the Hα line for the four different microturbulent velocity models. The model $\nu_3 = \nu_{F1} - \nu_{F2}$ is the one with the highest Hα intensity. The intensity decreases with the increase of the microturbulent velocity.

6.5 Conclusions

Comparing tables 6.2 and 6.4, we can see that the intensity in the Ly-β is lower than Lyα, probably because it is more difficult to populate level 3 as it requires more energy; therefore level 3 is normally less populated than level 2. In addition, level 3 can be de-excited and yield a transition either to level 2 (Hα), or to level 1 (Ly-β).
This is called a ‘branching’ between Ly-β and Hα.

For the Lyα and Hα lines, the intensity increases from the limb of the Sun to the centre of the disk; instead, the Ly-β line has a different behavior, the intensity increases from the centre to the edge of the Sun. We could expect to have higher intensities at the limb than at the centre because the variation of the line-of-sight is, in principle, intercepting more material when we look at the limb than at the centre of the disc, but actually we should take into account other factors like whether the line is in emission or in absorption, the temperature structure of the atmosphere, the height formation of the line or its optical thickness, etc.

The optically thick lines are not affected by the limb darkening effect: we expect to have a higher intensity at the limb than at the centre of the Sun because the variation of the line-of-sight is, in principle, intercepting more material when we look at the limb than at the centre of the disc. The Ly-β line is an optically thick line and for optically thick lines it is trickier and difficult to predict this effect.

We observed that the variation of intensity is related with the temperature structure of the model. The location of the transition region plays an important role: model \( T_2 \) with a low transition region yields the largest intensities. The variation of the temperature in the chromosphere is also related to the intensity, in fact the intensity associated to the model \( T_4 \) is lower than the one of the model \( T_1 \).

Concerning the variation of the microturbulent velocity at different heights, even if the results do not change significantly, the intensity decreases with the increase of the microturbulent velocity. The increase of the microturbulent velocity broadens the line profile, being broader for \( \nu_2 \) and narrower for \( \nu_1 \).

The microturbulent velocity plays a role in the equation of state when solving the hydrostatic equilibrium. When the code determines the hydrogen population densities at each position, the microturbulent velocity affects the shape of the spectral lines, depending on the actual transition, the optical thickness of the line, etc. meaning that at larger microturbulent velocities, the hydrogen densities will be lower. This fact could explain the decrease of the Ly-β intensity with the increase of the microturbulent velocity.
Chapter 7

A Radiative MagnetohydroDynamic Code applied to solar flares

7.1 Objective

I will describe the application of a radiative magnetohydrodynamic code to solar flares, in order to evaluate the response of the chromosphere to the deposition of thermal energy, which marks the optical lines and the continuum radiation observed during flares.

The aim of this chapter is to create a 1D plane-parallel atmospheric model which provides a detailed atmospheric diagnostic in order to compare it with the latest observations.

In previous works, several people as Abbott & Hawley (1999); Allred et al. (2005); 
Kašparová et al. (2009) have studied the dynamical response of the lower atmosphere to a flare energy input. Basing on their studies, I will use a radiative magnetohydrodynamic code from Carlsson & Stein (1992, 1997) to study the emission of the H\textalpha and Ly\alpha lines and compare them with the previous observations from Chapters 4 and 5.

7.2 The equation set

The code solves, in an adaptive grid (Dorfi & Drury 1987), the following equations:

- The statement of the conservation of mass (where there are no sources or sinks). Assuming a one-dimensional plane-parallel atmosphere, the plasma properties depend only on depth and the continuity equation reduces to:

\[ \frac{\partial \rho}{\partial t} + \frac{\partial \rho v}{\partial z} = 0 \]  

(7.1)

- The equation of momentum conservation:

\[ \frac{\partial (\rho v)}{\partial t} + \frac{\partial (\rho v^2)}{\partial z} + \frac{\partial p}{\partial z} = -\rho g \]  

(7.2)

- The equation of internal energy conservation:

\[ \frac{\partial (\rho e)}{\partial t} + \frac{\partial (\rho v e)}{\partial z} + (p + q_v) \frac{\partial v}{\partial z} + \frac{\partial}{\partial z} (F_c + F_r) = Q \]  

(7.3)
• The transfer equation:
  \[ \frac{dI_\nu}{dz} = j_\nu - \alpha_\nu I_\nu \]  
  (7.4)

• The charge conservation equation:
  \[ \sum_k n_{ik} I_{ik} + n_H n_e^{\text{met}}(T, n_e) = n_e \]  
  (7.5)

• The level population equations:
  \[ \frac{\partial n_{ik}}{\partial t} + \frac{\partial n_{ik} \nu}{\partial z} = \sum_{j \neq i}^N (n_{jk} P_{jik} - n_{ik} P_{ijk}) \]  
  (7.6)

• The atomic abundances:
  \[ \sum_i n_{ik} = A_k n_H \]  
  (7.7)

In these equations \( z, t, \rho, \nu, e \) and \( p \) have the usual definitions of height, time, mass density, velocity, internal energy per unit mass and gas pressure respectively. The \( q_\nu \) is the viscous pressure, and \( g \) is the acceleration of gravity. \( Q \) is any additional non-radiative heating required to give a specific stationary temperature distribution and \( F_c \) and \( F_r \) refer to the conductive and radiative fluxes. \( j_\nu \) is the emissivity and \( \alpha_\nu \), the extinction coefficient. \( n_{ik} \) is the population density of the energy level \( i \) of the element \( k \). \( I_{ik} \) is the ionization degree of energy level \( i \) of element \( k \). \( n_H \) is the total population number of hydrogen (neutral and ionised), \( n_e^{\text{met}}(T, n_e) \) is the number of electrons per hydrogen nucleus as a function of temperature and electron density. \( P_{ijk} \) is the probability of a transition from energy level \( i \) to energy level \( j \) in element \( k \). \( A_k \) is the abundance of element \( k \) on a scale where the abundances of hydrogen is 1. In Maltby & Leer (1990) there is more information about this set of equations.

The radiative transfer equation is solved using the linearization process proposed by Scharmer & Carlsson (1985), using a pentagonal lambda operator (Rybicki & Hummer 1991). The adaptive grid (Dorfi & Drury 1987) is used to manage with the steep gradients in the regions where the atomic level populations present sudden fast changes.

### 7.3 Atomic parameters

A six level hydrogen atom, a six level singly ionized calcium atom, a nine level helium atom and a four level singly ionized magnesium atom are treated in non-LTE. More information about the structure of these atoms can be found in Carlsson & Stein (2002). The atomic states and the line and continuum transitions are listed in Table 7.1 from Allred et al. (2005). The other atoms are included in the calculations as background continua in LTE using the Uppsala opacity package and applying the Saha-Boltzmann equations.

### 7.4 Pre-flare atmosphere

It is necessary to have a pre-flare atmosphere as a starting point for the flare simulations. In these simulations I will use the preflare atmosphere of Abbett & Hawley 1

\(^1\)0 is for a neutral element, 1 for a single ionized element, etc.
### 7.4. PRE-FLARE ATMOSPHERE

**Table 7.1: Bound-Bound transitions treated in non-LTE.**

<table>
<thead>
<tr>
<th>Atom</th>
<th>$\lambda_{ij}$ (Å)</th>
<th>Transition</th>
<th>Atom</th>
<th>$\lambda_{ij}$ (Å)</th>
<th>Transition</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>1215.70</td>
<td>$Ly\alpha$</td>
<td>Ca II</td>
<td>8662.16</td>
<td>$2d_3 \leftrightarrow 2p_1$</td>
</tr>
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<td>Ca II</td>
<td>8498.01</td>
<td>$2d_3 \leftrightarrow 2p_3$</td>
</tr>
<tr>
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<td>Ca II</td>
<td>8542.05</td>
<td>$2d_5 \leftrightarrow 2p_3$</td>
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<tr>
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<td>He $\alpha$</td>
<td>He I</td>
<td>601.42</td>
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<tr>
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<td>H $\beta$</td>
<td>He I</td>
<td>10830.29</td>
<td>$1s 2s^2 S_1 \leftrightarrow 1s 2p^3 P^o_4$</td>
</tr>
<tr>
<td>H I</td>
<td>4340.62</td>
<td>H $\gamma$</td>
<td>He I</td>
<td>584.35</td>
<td>$1s^2 1S_0 \leftrightarrow 1s 2p^1 P^o_1$</td>
</tr>
<tr>
<td>H I</td>
<td>18752.27</td>
<td>P $\alpha$</td>
<td>He I</td>
<td>20580.82</td>
<td>$1s 2s^1 S_0 \leftrightarrow 1s 2p^1 P^o_1$</td>
</tr>
<tr>
<td>H I</td>
<td>12818.86</td>
<td>P $\beta$</td>
<td>He II</td>
<td>303.79</td>
<td>$1s^2 S_{1/2} \leftrightarrow 2s^2 S_{1/2}$</td>
</tr>
<tr>
<td>H I</td>
<td>40513.47</td>
<td>Bo $\kappa$</td>
<td>He II</td>
<td>303.78</td>
<td>$1s^1 S_{1/2} \leftrightarrow 2p^2 P^o_2$</td>
</tr>
<tr>
<td>Ca II</td>
<td>3968.46</td>
<td>$2s \leftrightarrow 2p_1$</td>
<td>Mg II</td>
<td>2800.41</td>
<td>$2p^6 3s^2 S_{1/2} \leftrightarrow 2p^6 4p^2 P^o_{1/2}$</td>
</tr>
<tr>
<td>Ca II</td>
<td>3933.65</td>
<td>$2s \leftrightarrow 2p_3$</td>
<td>Mg II</td>
<td>2795.59</td>
<td>$2p^6 3s^2 S_{1/2} \leftrightarrow 2p^6 4p^2 P^o_{3/2}$</td>
</tr>
</tbody>
</table>

Figure 7.1: The initial temperature stratification of the PF2 atmosphere model of Abbett & Hawley (1999), compared with the Val3c and Mavn2 models of Vernazza et al. (1981) and Machado et al. (1980) for Quiet Sun and Bright Flares. Shown is the temperature (K) as a function of the column mass (g cm$^{-2}$) in logarithmic scale.

(1999). Fig. 7.1 shows the pre-flare temperature profile in comparison with the one of Vernazza et al. (1981) for the quiet sun, presented in the previous Chapter 6.

Fig. 7.2 shows the variation of the temperature, density and electron density with height for the pre-flare atmosphere used during the simulations.

Once a LTE solution is found for the atmosphere, it is used as the starting atmosphere to determine the non LTE populations of hydrogen. Thus, the atmospheric model will converge to a solution, which will be used as a background solution to find the non LTE populations densities of helium, calcium and magnesium atoms.

The atmosphere model is used to solve the radiation hydrodynamic equations of Section 7.2. As boundary conditions, the temperature at the apex of an initial coronal loop (see Fig. 7.4) is fixed at 10$^6$ K, and at the bottom, a non radiative heating is applied to fix the gradient of the temperature at the base of the photosphere.
7.5 External heatings

The RADYN code was initially developed by Carlsson & Stein (1992, 1997). In order to simulate the flare atmosphere, from the corona to the photosphere, Abbett & Hawley (1999) introduced a flux of energetic, accelerated electrons at the corona and a soft X-ray irradiation due to the flare heating. Later on, Allred et al. (2005) improved it by modifying the non-thermal electron beam input and adding EUV heating as well.

7.5.1 Electron beam heating

Abbett & Hawley (1999) calculated the nonthermal heating rate from Emslie (1978):

\[
Q_e(N) = \frac{\pi e^4}{\mu_0} \gamma(N) (\delta - 2) B_{xe} \left[ \frac{\delta}{2} \frac{1}{3} \right] \frac{F}{E} \left( \frac{N^*(N)}{N_c^*} \right)^{-\delta/2},
\]

where \( e \) is the fundamental unit of electric charge, \( x \) is the hydrogen ionization fraction as a function of column depth and \( \gamma = x \Lambda + (1 - x) \Lambda' \). \( \Lambda \) and \( \Lambda' \) are the Coulomb logarithms defined by Emslie (1978).

The values used for \( \Lambda \) and \( \Lambda' \) are those of Ricchiardelli (1982): \( \Lambda = 65.1 + 1.5 \ln E - 0.5 \ln n_h \), and \( \Lambda' = 25.1 + \ln E \) (the same values used in Hawley & Fisher (1994)). The electrons are injected at the coronal apex, and they propagate along magnetic field lines (assumed vertical and invariant at all heights), thus the pitch angle, \( \mu_0 \), is unity, \( F \) is the time dependent non-thermal electron energy flux that enters the magnetically confined loop at its apex. The quantity \( B_{xe}(\delta/2, 1/3) \) is the maximum column depth, given by \( N^*(N) = \int_{N_i}^{N} \gamma(N')/\Lambda dN' \), and \( N_c^* \) is the cutoff column depth for electrons in an ionized plasma \( (x = 1, \gamma = \Lambda) \). Details on the derivations of these expressions can be found in Emslie (1978), Emslie (1981) and Hawley & Fisher (1994) and the variation of the electron heating is shown in Figure 7.3. The resultant non-thermal heating is included in the equation of internal energy (7.3) as a source of external heating.

Later on, Allred et al. (2005) have modified the electron beam heating rate using hard X-ray spectra observations of a X-class solar flare that happened on 23th July 2002, taken with RHESSI (Holman et al. 2003), where the injected electrons are found to have a double power-law energy distribution.

Figure 7.2: The pre-flare atmosphere. (a): Temperature stratification; (b): Density stratification; (c): Electron density stratification.
7.5. EXTERNAL HEATINGS

\[
F_0(E_0) = \frac{\mathcal{F}(\delta_u - 2)(\delta_l - 2)}{E_c^2[(\delta_u - 2) - (E_B/E_C)^{2-\delta_i}(\delta_u - \delta_l)]}
\times \begin{cases} 
\left( \frac{E_0}{E_C} \right)^{-\delta_i} & \text{for } E_0 < E_B, \\
\left( \frac{E_B}{E_C} \right)^{\delta_u - \delta_l} \left( \frac{E_0}{E_C} \right)^{-\delta_u} & \text{for } E_0 > E_B, 
\end{cases}
\] (7.9)

where \( \mathcal{F} \) is the electron energy flux that enters the magnetic loop, \( E_c \) is the cutoff energy already defined in Section 2.7.2.1 and \( E_B \) is the break energy where the distribution shifts from spectral index \( \delta_l \) to \( \delta_u \). Using the Eq. 7.9, Allred et al. (2005) calculated the energy deposition rate as a function of column mass depth:

\[
Q_e(N) = \frac{2\pi e^4 \mathcal{F}}{2\mu_0 E_c^2} \cdot \gamma(N) 
\times \begin{cases} 
\left( \frac{N^*(N)}{N_c^*} \right)^{-\delta_i/2} B_{x_c} \left( \frac{\delta_l}{2}, \frac{1}{3} \right) - B_{x_B} \left( \frac{\delta_u}{2}, \frac{1}{3} \right) & \text{for } E_0 < E_B, \\
\left( \frac{N^*(N)}{N_c^*} \right)^{-\delta_u/2} \left( \frac{E_B}{E_c} \right)^{\delta_u - \delta_l} B_{x_s} \left( \frac{\delta_u}{2}, \frac{1}{3} \right) & \text{for } E_0 > E_B, 
\end{cases}
\] (7.10)

Fig. 7.3 shows the beam heating rate, changing in height. The \( Q_e(N) \) resulting from Eq. 7.11 penetrates deeper in the atmosphere than the one of the Eq. 7.8.

![Figure 7.3: Comparison of the electron beam heating rate in the pre-flare atmosphere calculated by Abbett & Hawley (1999) and Allred et al. (2005): the solid line shows the heating rate used in Eq. 7.11 and the dashed line shows the heating rate of Eq. 7.8. Image taken from Allred et al. (2005).](image)

7.5.2 X-ray heating

Abbett & Hawley (1999) calculated the X-ray heating rate following the geometry of the coronal loop shown in Fig. 7.4, getting the Eq. 7.11. It was included in the equation of internal energy (Eq. 7.9) as a source of additional heating (\( Q \)), ensuring
detailed energy balance at all depths in the atmosphere and at all times during the flare.

\[ Q_j = \int_V n_j (h\nu - X_j) \sigma_{j\nu} \frac{F_\nu}{h\nu} \, d\nu, \]  

(7.11)

where \( n_j \) is the number density for ion \( j \), \( X_j \) is the ionization potential, \( \sigma_{j\nu} \) is the photoinization cross section and \( F_\nu \) is the X-ray flux.

Allred et al. (2005) have modified the X-ray heating rate, including an extension of the atomic transitions, increasing the wavelength range between 1 and 2500 Å, adding EUV and UV heating as well. The XEUV heating rate is given by the Eq. 7.12:

\[ Q_{XEUV} = \int \frac{F_\nu}{h\nu} \left[ \sum_j n_j (h\nu - X_j) \sigma_{j\nu} \right] \, d\nu, \]  

(7.12)

where the notation is the same than at the Eq. 7.11.

Fig. 7.5 shows the X-ray heating rate, changing in height. The dashed line represents the soft X-ray heating, \( Q_j \), from Eq. 7.11 calculated by Abbett & Hawley (1999) and the solid one represents the X-ray and ultraviolet heating, from Eq. 7.12 calculated by Allred et al. (2005). Comparing both profiles, the one from Allred et al. (2005) is 10 times larger than the one of Abbett & Hawley (1999) at the peak (at other heights the difference is less important).

7.6 Flare Energetic Models

To investigate the properties of the lower atmosphere during the impulsive phase of a solar flare, I used the preflare atmosphere PF2 from Abbett & Hawley (1999) at the RADYN version of Allred et al. (2005). I considered a cutoff energy of \( E_c = 20 \)
Figure 7.5: Comparison of the XEUV heating rate from Allred et al. (2005) with the soft X-ray heating rate of Abbett & Hawley (1999). Image taken from Allred et al. (2005).

and three levels of non-thermal electron energy flux in order to differentiate weak \((F = 10^9)\), moderate \((F = 10^{10})\) and bright flares \((F = 10^{11})\) ergs cm\(^{-2}\) s\(^{-1}\) (hereafter called runs F9, F10 and F11 respectively).

The quiescent atmosphere is heated for 60 seconds in the F9, 7.00 seconds in the F10 and for a shorter 1.50 seconds burst in the F11 run, which is computationally much more problematic.

7.6.1 Weak Flares

Fig. 7.6 shows the variation of the temperature (first row), density (second row) and electron density (third row) with the height above \(\tau_{5000} = 1\) (Mm).

In the first row we can see how the beam of electrons heats the atmosphere, moving the transition region to deeper layers in time. We can also distinguish two distinct ‘phases’: the gentle evaporation phase, where there are two temperature peaks at the beginning (Figs. 7.6(a) and (b)) and only one later (Fig. 7.6(c)) which are due to the deposition of the energy by non-thermal electrons. The explosive phase (Fig. 7.6(d)) is characterized by the formation of large compression waves and shocks which propel mass upwards into the transition region and corona and downwards towards the temperature minimum.

Initially most of the energy is deposited in the upper chromosphere. The electron beam heats the atmosphere very fast and the transitions between excited states of hydrogen radiate the energy away. After 0.3 seconds, the non-thermal electrons have heated the atmosphere enough to ionize a significant part of hydrogen (as it is shown in the first column of the Fig. 7.7), releasing the energy away.

In Fig. 7.6(b) we can see that there is not enough neutral hydrogen in the atmosphere to radiate the energy and the increase of the temperature starts to ionize also the helium atom, giving rise to the second temperature peak.

This ‘gentle’ phase continues to move the transition region downwards and to

\(^4\)We choose this value as a lower range to not consider the thermal component of the spectrum as it is explained in Section 2.7.2.1 and as a higher range to model a flare not very energetic
heat the corona until the moment in which the atmosphere is not able to radiate the energy efficiently anymore.

After 58 seconds, the atmosphere is not able to radiate the energy anymore and the temperature starts to increase abruptly forming a bubble, visible in Fig. 7.6(d). The pressure increases forcing the material to move away and forming shock waves upwards through the corona and downwards into the chromosphere, as Fig. 7.8 demonstrates.

The simplified acoustic waves equation is given by Eq. 7.13:

$$\frac{\partial^2 P}{\partial x^2} - \frac{1}{v^2} \frac{\partial^2 P}{\partial t^2} = 0,$$

(7.13)

where $P$ is the acoustic pressure and $v$, the velocity. The solution of this equation is proportional to the quantity $v \times \rho^{0.5}$, as Edwin & Roberts (1982) shows. Fig. 7.8 represents the variation of the quantity $v \times \rho^{0.5}$, giving us information about the motion of the acoustic waves at different times. Acoustic waves are created mainly by the turbulent flow close below the surface and this graphic is giving us information about how the perturbation generated by the electron beam affects the motion of the plasma. The blue is interpreted as plasma moving downwards due to the radiation loses$^3$ and the yellow means that the beam has heated the atmosphere and the plasma is moving upwards.

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$^3$radiation loses means decrease of the temperature and as consequence, a motion of the plasma
7.6. FLARE ENERGETIC MODELS

Figure 7.7: Variation of the hydrogen population density at the ground level, $n=0$ (first row), first energy level, $n=1$ (second row) and third energy level, $n=2$ (third row) for F09. The solid line represents the atmospheric state at different times (0.3, 3, 20 and 60 seconds); the dotted line represents the structure of the initial state.

Figure 7.8: Variation of the velocity of the plasma multiplied by the square root of the density with the height in time for F09. The black color means downwards motion; the white, upwards motion and red means zero motion.

downwards.
Figure 7.9: Contribution of the different components of the energy to the total energy for F09: the heating due to the electron beam (light blue); the heating function (green line); the conductive heating (orange line); the radiative heating due to optically thick transitions calculated in detail (dark blue solid line); the radiative heating due to optically thin metal cooling (dark blue dashed line); the work done by the pressure (red solid line) and the viscous dissipation (red dashed line).

Fig. 7.9 shows how the different heating and cooling processes affect the total power (in erg s$^{-1}$ g$^{-1}$) at different times. In Fig. 7.9(a), at 0.3 seconds, the main contribution comes from the heating by the non-thermal electrons beam, contributing although the radiative term between 2.5 and 5 Mm. After 3 seconds, in Fig. 7.9(b), again beam heating is the main contribution to the total energy, but the radiative loses are becoming more important. The conductive term is the second main contribution to the total power, especially between 1.27 and 1.35 Mn. In Fig. 7.9(c), 20 seconds later, the beam of electrons has moved into a deeper level (the maximum of its power is situated at $\approx$ 1.5 Mm) but the distribution of the energy is similar than at 3 seconds. The same behavior is observed at 60 seconds, in Fig. 7.9(d), where the power associated to the electron beam is situated at 2.5 Mm. The viscous dissipation is negligible everywhere.

Fig. 7.10 gives us an idea about how the total power is partitioned between the thermal energy and the ionization energy at different times. In Fig. 7.10(a), at 0.3 seconds, the main contribution comes from the heating by the non-thermal electrons beam, contributing although the radiative term between 2.5 and 5 Mm. After 3 seconds, in Fig. 7.10(b), again beam heating is the main contribution to the total energy, but the radiative loses are becoming more important. The conductive term is the second main contribution to the total power, especially between 1.27 and 1.35 Mn. In Fig. 7.10(c), 20 seconds later, the beam of electrons has moved into a deeper level (the maximum of its power is situated at $\approx$ 1.5 Mm) but the distribution of the energy is similar than at 3 seconds. The same behavior is observed at 60 seconds, in Fig. 7.10(d), where the power associated to the electron beam is situated at 2.5 Mm. The viscous dissipation is negligible everywhere.

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4. Let’s remember from Section 1.2.1 that, for the quiet sun, the photosphere is the inner surface of the Sun, with $\sim$ 500 km thick; the chromosphere arrives to $\sim$ 2500 km and the corona starts $\sim$ from 3000 km.
7.6. FLARE ENERGETIC MODELS

seconds, the main contribution to the total power between 1.0 and 1.55 Mm is due to the ionization of the hydrogen atom; for deeper levels, the thermal energy is the main component, becoming more important at higher layers in time, i.e., in Fig. 7.10(c), after 20 seconds, the thermal component becomes important at 1.4 Mm. In all the four cases, the contribution of the ionization of the other elements to the total power is very small.

7.6.2 Moderate Flares

For moderate flares, the flux of the injected beam of non-thermal electrons is \( F = 10^{10} \text{ ergs cm}^{-2} \text{ s}^{-1} \) and as it is shown in Fig. 7.11, the response and evolution of the atmosphere follows a similar behaviour but in a shorter time.

Initially, most of the energy is deposited locally in the upper chromosphere. The electron beam heats the atmosphere, exciting the hydrogen atoms. As Fig. 7.11(a) shows, this energy is radiated away efficiently. After 0.3 seconds (first column of the Fig. 7.11), the hydrogen is being ionized (second row of Fig. 7.12) and the energy is not radiated effectively anymore, increasing the temperature and giving rise to the first peak observed in Fig. 7.11(a); the temperature continues increasing till the atmosphere enters in an equilibrium state in which the heat of the electron beam is balanced by the radiative loses. As a consequence of the increase of temperature, the other elements start to ionize as well at higher levels, creating the second peaks.
observed in Fig. 7.11(b) at 1.2 Mm. As time goes, the temperature increases and this ‘gentle’ phase continues for 6.2 seconds, until the moment in which the atmosphere is not able to radiate away the energy efficiently anymore. From this moment the ‘explosive’ phase starts: the temperature starts to increase suddenly creating a bubble (Fig. 7.11(d)); the pressure increases moving the plasma away from the point of the maximum heating, as Fig. 7.13 shows.

After ~ 6.2 seconds, when the temperature starts to suddenly increase, the Fig. 7.13 shows that there is some plasma moving upwards due to the increase of the pressure (yellow color) surrounded by a motion of the plasma downwards (black color) due to the radiation losses: the temperature of the plasma decrease and consequently it starts to move downwards.

Fig. 7.14 shows the different power terms and their contribution to the total energy at different times. In Fig. 7.14(a), at 0.3 seconds, the main contribution comes from the conductive heating, which is balanced with the radiative losses term due to the optically thin metal transitions calculated in detail. There are two main peaks at 1.15 Mm and 1.52 Mm, which are associated with the two temperature peaks observed in Fig. 7.11(a), corroborating that the temperature peak is because the atmosphere is not able to radiate the energy efficiently anymore. After 1.7 seconds, in Fig. 7.14(b), again the conductive heating is balanced by the radiative losses term due to the cooling of the optically thin metals; the two temperature peaks are present as well. In Fig. 7.14(c), 5 seconds later, the distribution of the energy is similar than at 2 seconds. After 7 seconds, in Fig. 7.14(d), during the ‘explosive’ phase, the main
7.6. FLARE ENERGETIC MODELS

Figure 7.12: Variation of the hydrogen population density at the ground level, n=0 (first row), first energy level, n=1 (second row) and third energy level, n=2 (third row) for F10. The solid line represents the atmospheric state at different times (0.3, 2.0, 5.0 and 7.0 seconds); the dotted line represents the structure of the initial state.

Figure 7.13: Variation of the velocity of the plasma multiplied by the square root of the density with the height in time for F10. The black color means downwards motion and the white, upwards motion.
contribution to the total energy comes from the heating of the electron beam. The viscous dissipation is negligible everywhere.

Figure 7.14: Contribution of the different components of the energy to the total one for F10 (see Fig. 7.9 for details).

Fig. 7.15 shows the thermal energy and the hydrogen ionization energy terms in order to study how they contribute to the total energy at different times. For the F10 flare model, the main contribution comes from thermal energy and ionization of metals. The hydrogen ionization energy is almost null, which agrees with the fact that the hydrogen is already ionized after 0.3 seconds.

7.6.3 Strong Flares

A significant higher level of thick target heating, corresponding to non-thermal energy flux of $F = 10^{11}$ ergs cm$^{-2}$ s$^{-1}$ (the F11 run), causes the atmosphere to respond in a different manner (See Fig. 7.16). Initially, the majority of the non-thermal electrons penetrate to a depth of 1.15 Mm, where they thermalize and heat the atmosphere to $\approx 10^4$ K within 0.3 seconds. As it is shown in Fig. 7.16, modest pressure gradients develop in response to this rapid temperature increase, and material begins to flow away from the point of maximum energy deposition. As the density increases, the transition region moves upward, and by 0.10 seconds the beam begins to deposit much of its energy at 1.1 Mm (shown in Fig. 7.20(b)) creating the weak, local temperature maximum at that height seen in Fig. 7.16(c) and increasing in time (Fig. 7.16(d)).
7.6. FLARE ENERGETIC MODELS

After $\sim 0.80$ seconds, there is an increase in the temperature (first row in Fig. 7.19) at the height where flare heating is strongest; a point which itself is not static and moves upwards towards the transition region, as Fig. 7.18 demonstrates. Fig. 7.19(d) shows that the power due to the optically thick transitions is no longer able to mitigate the influx of beam energy, being the contribution of the electron beam the main one to the total power. The mass flow increase and the point of maximum flare heating moves higher, inducing a rapid rise of the temperature as time goes on.

As it is shown in Fig. 7.17, the flare heating has increased the number of thermal collisions, populating the excited states of atoms in the region. After 1.50 seconds (fourth column in Fig. 7.17), there is a minimum of hydrogen density, associated with the temperature bubble seen in Fig. 7.16(d). The atmosphere is not able to radiate the energy efficiently anymore and the temperature increases suddenly.

From Fig. 7.18 we get information about the motion of the acoustic waves and we easily see that after $\sim 0.8$ seconds (when the temperature suddenly increases), there is some plasma moving upwards do to the increase of the pressure (yellow color) surrounded by a motion of the plasma downwards (black color) due to the radiation losses.

Fig. 7.19 shows how the different parameters affect the total power at different times. In Fig. 7.19(a), we see that the main contribution comes from the beam heating (light blue line), as in the previous flare models.

After 0.03 seconds (Fig. 7.19(a)), the radiative heating of optically thick transi-
Figure 7.16: Variation of the temperature (first row), density (second row) and electron density (third row) with the height above $\tau_{5000} = 1$ (Mm) for F11. The solid line represents the atmospheric state at different times (0.03, 0.10, 0.82 and 1.50 seconds); the dotted line represents the structure of the initial state.

In Fig. 7.19(b), again beam heating is the main contribution to the total power, but the radiative losses become the main source of energy losses, changing in 0.07 seconds from a positive value to a negative value between 1.2 and 1.5 Mm. Between 1.35 and 1.52 Mm the conductive term contributes as a second main source to the total power. 0.82 seconds later, in Fig. 7.19(c), the terms associated to the radiative heating and the conductive heating moved to upper layers. After 1.50 seconds (Fig. 7.19(d)), the conductive heating term has significantly decreased and the contribution of the radiative losses is located in a narrower height range.

Fig. 7.20 gives us an idea about how the thermal energy and the ionization energy contribute to the total energy at different times.

In Fig. 7.20(a), at 0.3 seconds, the main contribution to the total energy between 1.05 and 1.2 Mm is due to the ionization of the hydrogen atom; after 1.56 Mm, the thermal energy is the main component, becoming more important at higher layers in time. After 0.1 seconds, the power (in erg s$^{-1}$ g$^{-1}$) due to the beam of electrons increases in intensity, but the total power does not correspond with the sum of all the contribution as in the other cases; this could be explained with the fact that for this flare model, at this time the temperature increases suddenly, without giving time to the atmosphere to react to the huge temperature increase. In all four cases, the contribution of the ionization of the other elements to the total energy is very small.
7.7. RESULTS

Figure 7.17: Variation of the hydrogen population density at the ground level, \( n=0 \) (first row), first energy level, \( n=1 \) (second row) and third energy level, \( n=2 \) (third row) for F11. The solid line represents the atmospheric state at different times (0.03, 0.10, 0.82 and 1.50 seconds); the dotted line represents the structure of the initial state.

Figure 7.18: Variation of the velocity of the plasma multiplied by the square root of the density with the height in time for F11. The black color means downwards motion and the white, upwards motion.

7.7 Results

In order to compare the results obtained from the model with the observed flares in Ly\( \alpha \) and H\( \alpha \) from Chapters 4 and 5, in the next Sections 7.7.1 and 7.7.2, we will
study the evolution and behavior of both Ly\(\alpha\) and H\(\alpha\) lines using the different flare atmosphere models.

7.7.1 Ly-\(\alpha\) emission

As it is explained in Chapter 4, from the observations of both solar flares in Ly\(\alpha\) using the TRACE 1216 Å channel, we were able to estimate the total intensity in Ly\(\alpha\) at the flare footpoints, being \((1 \times 10^6 - 1.8 \times 10^7)\) erg \(s^{-1}cm^{-2}sr^{-1}\), as it is shown in Table 4.10.

Calculating the intensity of the Ly\(\alpha\) line along the profile over an interval of \(\Delta \lambda = 5\) Å around line centre at different times, Fig. 7.21 shows the light curves for the three different flare models. The strength of the non-thermal electron beam flux (\(F\)) weakly affects the Ly\(\alpha\) peak intensity in time. The peak intensity is only less than a factor 2 higher for F11 than for F9. However the evolution of the light curve of the different flare models has different behaviors - For F09 (Fig. 7.21(a)), the intensity rises sharply in just 3 seconds, it then continues increasing for 14 seconds, and after that it decreases in time - keeping always the same order of magnitude. Instead for moderate flares, as Fig. 7.21(b) shows, the Ly\(\alpha\) intensity increases suddenly in 0.3 seconds; after that it remains almost constant for almost 6 seconds and during the last second associated to the ‘explosive’ phase, it continues increasing in time, but more slowly. For bright flares, (Fig. 7.21(c)), the behavior
7.7. RESULTS

(a) 0.03 seconds  
(b) 0.10 seconds  
(c) 0.82 seconds  
(d) 1.50 seconds

Figure 7.20: Contribution from the thermal and ionization energy to the total energy for F11.

of the light curve is similar to moderate flares: the intensity increases in time; after 0.7 seconds, there is an abrupt emission peak, most probably due to the abrupt temperature rise, with a duration of $\sim 0.05$ seconds and after that the intensity continues to rise with numerous spikes, which are due to the numerical calculations because the atmosphere is not stable anymore.

Figure 7.21: Variation of the Ly $\alpha$ intensity in time for the three flare models for a wavelength interval of $\Delta \lambda = 5$ Å around the line centre.
CHAPTER 7. A RADIATION MHD CODE

7.7.1.1 Faint Flares

Fig. 7.22 shows the Ly$\alpha$ line profile at different times for faint flares. Initially, the profile is a typical Ly$\alpha$ profile with two emission peaks. The line becomes broader in time; after 3 seconds, the blueshifted peak becomes stronger and the redshifted weaker; after 20 seconds, the redshifted peak is out of the wavelength range.

Figure 7.22: Variation of the Ly$\alpha$ intensity in time for the F09 flare model for a wavelength interval of $\Delta \lambda = 62$ Å.

To better understand how the Ly$\alpha$ line is formed, we studied the different contributions to the total intensity and how they vary. Taking into account the formal solution of the transfer equation, which is given by Eq. 7.14, the integrand term is called contribution function, $C_i$, it represents the fraction of the emergent intensity emanating from height $z$ at frequency $\nu$ and it is divided into three different terms:

$$I_\nu(0) = \frac{1}{\mu} \int_{\tau_\nu} S_\nu e^{-\tau_\nu/\mu} d\tau_\nu = \frac{1}{\mu} \int z S_\nu \frac{\chi_\nu}{\tau_\nu} e^{-\tau_\nu/\nu} dz,$$

(7.14)

- $\frac{\chi_\nu}{\tau_\nu}$, where $\chi_\nu$ is the extinction coefficient and $\tau_\nu$ is optical depth.

- $S_\nu$, which is called the source function and it is defined as $\frac{j_\nu}{\chi_\nu}$, where $j_\nu$ is the emissivity.

- $\tau_\nu e^{-\tau_\nu/\nu}$, which is the attenuation caused by the optical thickness.

Fig. 7.23 represents $\frac{\chi_\nu}{\tau_\nu}$ on the top left, $\tau_\nu e^{-\tau_\nu/\nu}$ on the top right, $S_\nu$ on the bottom left and $C_i$ on the bottom right. The color scale is represented on the right, where the dark is the minimum of intensity and the light is the maximum of intensity.

The main contribution to the $C_i$ at the chromosphere ($\sim 1.5$ Mm) is from the $\tau_\nu e^{-\tau_\nu/\nu}$ component; this term tells us at which level the attenuation factor is minimum, yielding a maximum of intensity in Ly$\alpha$ between 1.25 and 1.55 Mm. The
7.7. RESULTS

Figure 7.23: Different components of the intensity contribution for Lyα after 3.0 seconds of flare heating for F09. The solid line indicates the emission at \( \tau_\nu = 1 \) and the dashed line is the plasma velocity. The four different panels are the different terms of the Eq. (7.14): \( \frac{\chi_\nu}{\tau_\nu} \) on the top left, \( \tau_\nu e^{-\tau_\nu/\nu} \) on the top right, \( S_\nu \) on the bottom left and \( C_i \) on the bottom right.

maximum of \( \frac{\chi_\nu}{\tau_\nu} \) is at 1.6 Mm, but is balanced by the minimum of the source function, \( S_\nu \), at this layer. From the representation of the intensity distribution of the contribution function, \( C_i \), we see that the Lyα line is a very broad line: the photons associated to the wings of the line comes from 1.3 Mm and the ones of the center of the line, from 1.55 Mm; there are 0.25 Mm of distance between the two emission layers for the Lyα line.

7.7.1.2 Moderate Flares

In Fig. (7.24) we can study the evolution of the Lyα line profile in time. The evolution of the line profile follows the same behavior than for faint flares: the blueshifted peak becomes stronger and the redshifted weaker. The intensity of the blueshifted peak increases in time, as it is shown also in Fig. (7.21) b).

Fig. (7.25) shows the different terms associated to the contribution function \( C_i \) of the Lyα line after heating the atmosphere with the electron beam for 2 seconds. In this case, the distribution of the function is very similar to the one of weak flares at 3 seconds: the Lyα maximum intensity is found between 1.15 and 1.45 Mm. The maximum of \( \frac{\chi_\nu}{\tau_\nu} \) is at 1.05 Mm and 1.50 Mm, but this second peak it is balanced by the minimum of the source function, \( S_\nu \), at this layer.

7.7.1.3 Bright Flares

Fig. (7.26) shows the Lyα line profile at different times for the F11 flare model. The evolution of the line profile follows the same behavior as the previous flare models: the line becomes broader in time; the blueshifted peak becomes stronger and the
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Figure 7.24: Variation of the Lyα intensity in time for the F10 flare model for a wavelength interval of $\Delta \lambda = 62 \, \text{Å}$.

Figure 7.25: Different components of the intensity contribution for Lyα after 2.0 seconds of flare heating for F10. See Fig. 7.23 for other details.

redshifted weaker. After 1.50 seconds, the redshifted peak is out of the wavelength range.

Fig. 7.27 represents the different terms associated to the contribution function $C_i$ of the Lyα line after heating the atmosphere with a flux of non-thermal electron
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(a) $t=0.03$ seconds

(b) $t=0.10$ seconds

(c) $t=0.82$ seconds

(d) $t=1.50$ seconds

Figure 7.26: Variation of the Ly$\alpha$ intensity in time for the F11 flare model for a wavelength interval of $\Delta \lambda =62$ $\AA$.

beam of $F = 10^{11}$ ergs cm$^{-2}$ s$^{-1}$ for 0.82 seconds. The maximum contribution of the $\frac{S_{\nu}}{\tau_{\nu}}$ term comes from 1.0 Mm; the source function, $S_{\nu}$, has its maximum between 1.1 and 1.25 Mm and the attenuation factor, $\tau_{\nu}e^{-\tau_{\nu}/\nu}$, contributes to $C_{i}$ between 1.0 and 1.3 Mm. Multiplying all the terms in order to get $C_{i}$, the resulting contribution function is maximum at the chromosphere, between 1.25 and 1.30 Mm. Therefore we see that for bright flares, the core and the wings of the Ly$\alpha$ are formed over a narrower region than for weak and moderate flares.

7.7.2 H-$\alpha$ emission

In Chapter 5 we described the evolution of a C7.3 solar flare, studying its H$\alpha$ light curve and variation of the H$\alpha$ line profile during the end of the impulsive phase. The RADYN code gives the opportunity to study how the beam heating affects the H$\alpha$ intensity and to compare the simulated results with the observations.

Fig. 7.28 shows the H$\alpha$ light curve during the flare for each run over a wavelength range of $\Delta \lambda =5$ $\AA$. The H$\alpha$ total integrated intensity shows a different evolution in time for each kind of flare, and the peak intensity varies by less than a factor of 2 between F09 and F11. For faint flares (Fig. 7.28(a)) the H$\alpha$ intensity increases sharply in just 3 seconds; after that remains constant for another 5 seconds; and then the intensity starts to decrease in time. For moderate flares (Fig. 7.28(b)), the intensity increases in 0.3 seconds and after that it remains constant. Finally, the H$\alpha$ intensity for bright flares increases abruptly for the first 0.05 seconds, continues increasing in time slowly, at 0.7 seconds occurs a peak; and later on it continues increasing keeping always the same order of magnitude.
7.7.2.1 Weak flares

The Hα profile is shown in Fig. 7.29(a) at different times. Initially, the Hα line is in absorption, as Canfield et al. (1990) observed. After 0.3 seconds there is a double peak emission in the wings of the line, the blueshifted peak being stronger; as time goes, the Hα intensity then decreases as the light curve from Fig. 7.28 demonstrates.

Fig. 7.30 represents the different terms associated to the contribution function $C_i$ of the Hα line after heating the atmosphere with a flux of non-thermal electron beam of $F = 10^9$ ergs cm$^{-2}$ s$^{-1}$ for 3 seconds. The maximum of the contribution factor is situated at 1.25 Mm for the wings of the line and at 1.35 Mm for the center of the Hα line, which verifies that the Hα line is formed in the chromosphere and that the photons from the center of the line are coming from higher layers than the ones from the wings of the line.

Figure 7.28: Variation of the Hα intensity in time for the three flare models for a width of $\Delta \lambda = 5$ Å.
7.7. RESULTS

(a) F09
(b) F10
(c) F11

Figure 7.29: Hα line profile for a width of $\Delta \lambda = 500$ Å at different times for the three different flare models. (a): Faint flares; (b): Moderate Flares; (c): Bright Flares.

Figure 7.30: Different components of the intensity contribution for Hα after 3.0 seconds of flare heating for F09. The solid line indicates the emission at $\tau_\nu = 1$ and the dashed line is the plasma velocity. The four different panels are the different terms of the Eq. 7.14: $\chi_\nu$ on the top left, $\tau_\nu e^{-\tau_\nu/\nu}$ on the top right, $S_\nu$ on the bottom left and $C_i$ on the bottom right.

7.7.2.2 Moderate flares

In Fig. 7.29(b) we can observe the Hα line profile at different times and how it evolves in time. After 0.3 seconds, the Hα line becomes a symmetric emission line with two emission peaks in the wings; as time goes, the blueshifted peak becomes stronger and the redshifted one weaker, as we observed for the Lyα line.

For moderate flares, the contribution function associated to the Hα line after heating the atmosphere with a flux of non-thermal electron beam of $\mathcal{F} = 10^{10}$ ergs cm$^{-2}$ s$^{-1}$ for 2 seconds is shown in Fig. 7.31. From the bottom right panel we can see that the formation region of the Hα line spans about 0.10 Mm, much narrower than the Lyα line. The maximum of the contribution function comes from 1.06 Mm
and it is associated with the wings of the Hα line. The intensity at the center of the line is affected by a blue shift, which can be interpreted as an upward motion of plasma through the chromosphere.

### 7.7.2.3 Bright flares

Fig. 7.29(c) shows the Hα line profile at different times for the F11 flare model. After 0.03 seconds, two emission peaks appear in the wings of the line, which become stronger in time, as opposed to the Hα line profile for F09. The line is now symmetric, with some fluctuations in its intensity and line profile.

The contribution function associated to the Hα line after heating the atmosphere with a flux of non-thermal electron beam of \( F = 10^{11} \) erg s\(^{-1} \) cm\(^{-2} \) for 0.82 seconds is shown in Fig. 7.32. In this case, the dashed line, associated to the velocity of the plasma is vertical, meaning that the motion is null at this level of the atmosphere, creating a symmetric Hα profile, as opposed to the other flare models.

### 7.8 Conclusions

Comparing the Lyα value estimated from TRACE images with the theoretical one obtained from the RADYN code for the three different flare models, both are comparable even if the simulated ones are slightly higher: in RADYN, the scattering is assumed to be isotropic and incoherent and the energy flux of the non-thermal electron beam is constant in time. These facts may explain the differences.

Table 7.2 shows the comparison between the observed Lyα intensity estimated for the two flares studied in Chapter 4, the Hα intensity of the flare studied in Chapter 5, the integrated intensity resulted from the radiative transfer code from Chapter 6 using the model \( T_3 \) and for a range of 2 Å for Lyα and 0.5 Å for Hα and

![Figure 7.31](image-url)  
Figure 7.31: Different components of the intensity contribution for Hα after 2.0 seconds of flare heating for F10. See Fig. 7.30 for other details.
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Figure 7.32: Different components of the intensity contribution for H\(\alpha\) after 0.82 seconds of flare heating for F11. See Fig. 7.30 for other details.

The integrated intensity resulted from the radiative MHD code from this Chapter for the three different flare models, for a wavelength interval of 10 Å for Ly\(\alpha\) and 20 Å for H\(\alpha\).

<table>
<thead>
<tr>
<th></th>
<th>Ly(\alpha) Intensity (erg s(^{-1}) cm(^{-2}) sr(^{-1}))</th>
<th>H(\alpha) Intensity</th>
</tr>
</thead>
<tbody>
<tr>
<td>Observations</td>
<td>(1 \times 10^6)-(1.8 \times 10^7)</td>
<td>(2 - 3.5 \times 10^6)</td>
</tr>
<tr>
<td>Radiative Transfer (T_3)</td>
<td>(3.2 \times 10^6)</td>
<td>(2.0 \times 10^7)</td>
</tr>
<tr>
<td>RADYN F09 (t=3 s)</td>
<td>(2.2 \times 10^6)</td>
<td>(7.7 \times 10^5)</td>
</tr>
<tr>
<td>RADYN F10 (t=2 s)</td>
<td>(7.2 \times 10^7)</td>
<td>(1.2 \times 10^7)</td>
</tr>
<tr>
<td>RADYN F11 (t=0.82 s)</td>
<td>(1.0 \times 10^8)</td>
<td>(1.5 \times 10^7)</td>
</tr>
</tbody>
</table>

Table 7.2: Comparison of the Ly\(\alpha\) and H\(\alpha\) observed intensities with the ones from the radiative transfer and RADYN codes.

The integrated intensity in the H\(\alpha\) line obtained from the observations is lower than the simulated ones. The behaviour of the two light curves is different (compare Figs. 7.28 with Fig. 5.11): this might be explained by the time variation of the beam flux.

The Ly\(\alpha\) intensity is affected by the flare model, being two orders of magnitude difference between bright and faint flare models, instead the H\(\alpha\) intensity is still more than one order of magnitude between F09 and F11.

Even if the results from RADYN match with the observations, the H\(\alpha\) intensity is better fitted than the Ly-\(\alpha\).
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Conclusions and Future Work

8.1 Conclusions

We focused on the behavior of the Lyα and Hα intensities because the study of both Lyα and Hα flares allows us to study the emission of solar flares at the chromosphere.

The study of the two solar flares in Lyα presents the first examination of a flare in TRACE 1216 Å at high spatial and temporal resolution. From the observations we concluded that the Lyα emission is well co-spacially related with the X-ray emission at the flare footpoints; the Lyα power is less than 10% of the power inferred for the electrons, under the assumption of a collisional thick target chromosphere, which means that it can be provided by the electrons which produce the observed X-ray radiation.

Concerning the flare observed on 08 September 1999, this event was associated with the eruption of a small filament, which was heated and erupted just before the beginning of the flare, emitting UV radiation from the chromosphere.

The 28 February 1999 flare was also associated with the destabilization of a filament: an initial filament destabilization acts as a trigger for a two-ribbons flare and fifteen minutes later it causes the activation of a second filament associated to the same active region. The filament develops in brightness giving rise to a CME.

The Lyα power estimated for the M1.4 flare observed on 08 September 1999 is $2.4 \times 10^{25} \text{ erg s}^{-1}$ and $8.1 \times 10^{25} \text{ erg s}^{-1}$ for the M6.6 flare observed on 28 February 1999: the more energetic flare according to the GOES X-ray classification is also the more energetic in Lyα.

On 03 July 2002, a C7.3 class flare was observed in Hα, showing a filament that was destabilized. The Hα intensity calculated for the C7.3 flare is $2 - 3.5 \times 10^6 \text{ erg s}^{-1} \text{ cm}^{-2} \text{ sr}^{-1}$.

Using the radiative transfer code, we studied how the temperature and micro-turbulence velocity and the variations of the solar center - limb affect the Lyα, Ly-β and Hα lines profiles and their intensities, in order to better understand how the chromospheric emission is affected by these parameters. The Lyα and Hα intensities increase from the limb to the centre of the Sun; instead, the Ly-β intensity has the opposite behavior, most probably because some parameters like the height formation of the line or the optical thickness affect in a different way this line.

Using different temperature atmospheric models, we could also study how the location of the transition region affects the intensities: a lower transition region in height involves a larger intensity in all the three lines (Lyα, Ly-β and Hα). The variation of the temperature in the chromosphere is also related to the intensity: a lower temperature of the chromosphere implies a lower intensity of the chromospheric
The increase of the microturbulent velocity implies a decrease of the chromospheric intensities, broadening the line profiles.

Comparing the resulting intensities from the radiative transfer code with the observed ones we can observe that for the Ly$\alpha$ intensities the results are the same order of magnitude than the observations for all the flare models. Instead, for the H$\alpha$ line, the resulted intensity is higher than the observations, most probably because the observed H$\alpha$ flare is a C7.3 class, not very energetic, instead the Ly$\alpha$ flares are classified according to GOES as M1.4 and M6.6, which are stronger than a C7.3 class flare.

From the RADYN code, taking into account the dynamical response of the solar chromospheres to the energy injected in the form of non-thermal electrons during solar flares, we studied the flare energy transport and radiation production in the chromosphere as well as the H$\alpha$ and Ly$\alpha$ emission.

We observed that the Ly$\alpha$ intensity is affected by the flux of the initial beam of electrons injected at the top of the loop, while the H$\alpha$ intensity is less affected by the flare model.

The integrated intensity in the H$\alpha$ line obtained from the observations is slightly lower than the simulated ones; the behavior of the simulated and observed light curves is different, most probably due to the time variation of the beam flux.

### 8.2 Future Work

As well as its intrinsic interest, we have been in fact motivated to perform this analysis by forthcoming instrumentation. The Extreme Ultraviolet Imaging Telescope (see [http://www.sidc.be/publications/docs/EUI-AthensPaper-20061120.pdf](http://www.sidc.be/publications/docs/EUI-AthensPaper-20061120.pdf) from Hochez et al. (2007)) which has been selected for ESA’s Solar Orbiter mission is intended to carry out Ly$\alpha$ imaging, as will the LYOT imager on the proposed France-China SMiSE satellite (Vial et al. 2007) currently under study. These Ly$\alpha$ instruments will provide an imaging overview of the flare evolution, both the flare ribbons/footpoints and the initial phases of the filament ejection, and will be particularly valuable scientifically when carried out in conjunction with hard X-ray imaging.

Using the Radiative Magnetohydrodynamic code, we can carry out the following investigations to see if the simulations are able to explain some observational effects:

1. EUV and soft X-ray imaging shows rapidly-varying brightenings (Hudson et al. 1994; Fletcher & Hudson 2001) implying fast heating and cooling. EIS and RHESSI (Milligan & Dennis 2009) also show that some hot material (up to 2 MK) is downflowing, in stark contrast to common evaporation models. With the RMHD code, by using a beam heating we are able to simulate the heating of the plasma in a short time scale and better understand its behavior.

2. Strong chromospheric lines are formed under non local thermodynamic equilibrium conditions, like H$\alpha$ and Ly$\alpha$. These lines can be influenced by non-thermal processes. Understanding the line formation is crucial for the correct interpretation of the observations; we can carry out detailed modeling of non-LTE effects on the line shape and evolution (e.g. Ding & Fang (2001); Berlicki (2007); Kašparová et al. (2007)).
8.2. FUTURE WORK

At the same time, we can work on the RADYN code taking into account the following improvements:

- Adding a proton beam and changing the energy deposition.
- Modelling a gradual phase, turning off the electron beam and keeping the atmosphere stable, in order not to cool it too fast.
- Adding a realistic energy flux of the non-thermal electron beam and the cut-off energy, from the RHESSI observations, changing in time.
- Changing the electron beam deposition, using the Fokker-Planck energy deposition analysis, improving the one used currently on the code.
- Including collisional excitations from the ground state, which most probably will affect to the resulted density populations.

The next generation of instruments (e.g. the IRSI spectrometer under development at LMSAL) promises chromospheric observations as a main focus. Meanwhile, the numerical techniques are now at the level of sophistication necessary to model the complexity of this most challenging layer of the atmosphere of the Sun. It is with these facts in mind that we want to return to the chromosphere.
It’s been a long way to write it, and it’s been great fun. This would have not happened without many of you. It sounds like a cliché but it is a very true one.

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