ON REDSHIFTS AND BLUESHIFTS IN THE TRANSITION REGION AND CORONA

V. H. Hansteen$^{1,2,3,4}$, H. Hara$^2$, B. De Pontieu$^3$, and M. Carlsson$^{1,4}$

$^1$ Institute of Theoretical Astrophysics, University of Oslo, P.O. Box 1029 Blindern, N-0315 Oslo, Norway; viggoh@astro.uio.no, mats.carlsson@astro.uio.no
$^2$ National Astronomical Observatory of Japan, 2–21–1 Osawa, Mitaka, Tokyo 181–8588, Japan; hirohisa.hara@nao.ac.jp
$^3$ Lockheed Martin Solar & Astrophysics Lab, Org. ADBS, Bldg. 252, 3251 Hanover Street Palo Alto, CA 94304, USA; bdp@lmsal.com
$^4$ Center of Mathematics for Applications, University of Oslo, P.O. Box 1053 Blindern, N-0316 Oslo, Norway

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ABSTRACT

Emission lines formed in the transition region (TR) of the Sun have long been known to show pervasive redshifts. Despite a variety of proposed explanations, these TR downflows (and the slight upflows in the low corona) remain poorly understood. We present results from comprehensive three-dimensional MHD models that span the upper convection zone up to the corona, 15 Mm above the photosphere. The TR and coronal heating in these models is caused by the stressing of the magnetic field by photospheric and convection “zone dynamics,” but also in some models by the injection of emerging magnetic flux. We show that rapid, episodic heating, at low heights of the upper chromospheric plasma to coronal temperatures naturally produces downflows in TR lines, and slight upflows in low coronal lines, with similar amplitudes to those observed with EUV/UV spectrographs. We find that TR redshifts naturally arise in episodically heated models where the average volumetric heating scale height lies between that of the chromospheric pressure scale height of 200 km and the coronal scale height of 50 Mm.

Key words: Sun: chromosphere – Sun: corona – Sun: transition region – Sun: UV radiation

Online-only material: animations

1. INTRODUCTION

The transition region (TR) acts as a conduit for all of the non-thermal energy that powers the outer solar atmosphere. However, the nature of the TR remains poorly understood (Mariska 1992), with observations of the upper chromosphere and of the TR pointing to a number of mysteries that have yet to be adequately explained. Chief among these are the rise in the differential emission measure (DEM) at temperatures below $10^5$ K (Mariska 1992) and the ubiquitous average redshift found in TR lines (e.g., Doschek et al. 1976; Lites et al. 1976; Dere et al. 1984; Achour et al. 1995; Brekke et al. 1997; Chae et al. 1998; Peter & Judge 1999).

A variety of models have been proposed to explain the average redshift of the TR emission lines. For example, Athay & Holzer (1982) and Athay (1984) conjectured that redshifts were caused by the return of previously heated spicular material. On the other hand, Hansteen (1993) argued that nanoflare induced waves generated in the corona could produce a net redshift in TR lines.

Recently, Peter et al. (2006) found that three-dimensional (3D) numerical models that span the photosphere to corona produce redshifts in TR lines and may also reproduce the observed DEM. In these models, coronal heating is due to the Joule dissipation of currents that are produced as the magnetic field is stressed and braided by photospheric motions. However, the ultimate source of the redshifts is not clear from these simulations. More recent work on 3D models proposes that these downflows are signatures of cooling plasma that drains from the coronal volume after loops are disconnected from the reconnection site that drives the initial heating (Zacharias et al. 2009). Draining plasma is also the focus of some recent one-dimensional (1D) models (Bradshaw & Cargill 2010).

In contrast, Spadaro et al. (2006) used 1D models of coronal loops to show that redshifts and the observed DEM at temperatures below $10^5$ K could be the result of transient heating near the loop footpoints.

In this work, we extend the work of Peter et al. (2006) and Spadaro et al. (2006) and show that TR redshifts are naturally produced in episodically heated 3D models in which the average volumetric heating scale height lies between that of the chromospheric pressure scale height of 200 km and the coronal scale height of 50 Mm. In order to do so, we briefly present results from 3D MHD models spanning the upper convection zone up to the corona (Section 2). The TR and coronal heating in these models is caused by the stressing of the magnetic field by photospheric and convection zone dynamics, but also in some models by the injection of emerging magnetic flux. In Section 3, we describe the characteristics of the simulated line profiles constructed on the basis of these models, and in Section 4, the heating mechanism that arises in our model is described. We finish the paper with a discussion of how this heating leads to TR redshifts that are similar to those observed and suggest other comparisons between models and observations.

2. 3D SIMULATIONS OF THE SOLAR ATMOSPHERE

We solve the MHD equations using the BIFROST code (B. V. Gudiksen et al. 2010, in preparation) as well as the older, but essentially identical Oslo Stagger Code (Hansteen et al. 2007). These codes employ a high-order finite difference scheme and include high-order artificial viscosity and resistivity in order to maintain numerical stability on a grid of finite resolution. These terms are also the source of magnetic and viscous heating. Non-gray, optically thick radiative losses including the effects of scattering in the chromosphere are included (Skartlien 2000), as are optically thick radiative losses in the chromosphere, optically thin radiative losses in the upper chromosphere and corona and conduction along the field lines in the corona. Further details of the codes and the equations solved can also be found in Martínez-Sykora et al. (2008, 2009).
2.1. Simulation Description

In this paper, we will use results from four 3D simulations of the solar atmosphere. These models span the region from $z = -1.4$ Mm (we use the standard convention of setting the zero point for our height scale to the average height where the optical depth at 500 nm is unity) to $z = 14.4$ Mm and thus include the upper convection zone, the photosphere, chromosphere, TR, and the (lower) corona. Photospheric and lower chromospheric radiative losses are balanced by an inflowing heat flux at the bottom boundary. The chromosphere and regions above are heated by the dissipation of acoustic shock waves that are generated in the convection zone and photosphere, but perhaps more importantly by the dissipation of the Poynting flux that is injected into the upper atmosphere. This Poynting flux arises as a result of the braiding of magnetic field lines as well as the injection of new “emerging” magnetic flux at the lower boundary. These processes have been described earlier in the context of numerical simulations of the outer solar atmosphere (Gudiksen & Nordlund 2005; Hansteen 2004; Hansteen & Gudiksen 2005; Hansteen et al. 2007) and also including flux emergence (Martnez-Sykora et al. 2008, 2009).

Three of the models discussed are comprised of $512 \times 512 \times 325$ points spanning a region of $16 \times 16 \times 15.5$ Mm$^3$. These models differ in their total unsigned magnetic flux. The weak field model (henceforth named “A1”) has $\langle |B_z| \rangle \approx 20$ G in the photosphere. The stronger field model (named “A2”) has $\langle |B_z| \rangle \approx 100$ G. In addition, in this stronger field model a flux tube is injected at the bottom boundary that emerges through the photosphere during the course of the simulation. Finally, the “A3” model is initially identical to A2 with $\langle |B_z| \rangle \approx 100$ G, but in this model very little flux is injected at the bottom boundary. All three models have been run for roughly 1 hr solar time.

The fourth model, “B1,” has lower resolution—$256 \times 128 \times 160$ points—and spans a smaller solar region of $16 \times 8 \times 15.5$ Mm$^3$. The unsigned magnetic field is of the same order, but slightly weaker $\langle |B_z| \rangle \approx 75$ G, than in the strong field models, A1 and A2, mentioned above. This low-resolution model also includes an injected emerging flux sheet and has been run for 1 hr solar time.

Typical for these models is a temperature structure in the TR and corona similar to that indicated by Figure 1 which shows the average, maximum and minimum temperatures as a function of height in the A2 model at $t = 2700$ s: the average temperature begins to rise some 0.7 Mm above the photosphere and has already reached 1.0 MK at heights of 3–4 Mm. Above this height, the average temperature continues to rise to some 2 MK at 10 Mm after which the temperature drops slightly. However, the structure of the atmosphere is far from horizontally homogeneous. This is illustrated by the plots of the minimum and maximum temperatures in Figure 1. We find very hot plasma (> 2 MK) already at heights of 1.5 Mm, while at the same time we find cool gas with temperatures < 10 kK even at heights up to 6 Mm. The strong field models, A2, A3, and B1, are all similar in this respect, while the weak field, A1, model remains cooler throughout the corona with average temperatures of order $5 \times 10^5$ K or below and maximum temperatures below $6 \times 10^5$ K.

The time evolution of the average coronal temperature can be illustrated by considering the average temperature at the upper boundary of our simulations (14.4 Mm) as functions of time as shown in Figure 2. Run A1 is relatively cool, starting with an average temperature of only 200 kK which rises to some 500 kK during the first 20 minutes of the simulation, very slowly cooling thereafter. Runs A2 and B1 both have higher pre-existing magnetic fields than run A1 and, in addition, have additional flux injected at the lower boundary. Hence, both models are initially hotter—of order 1 MK—and the average temperature rises further some 20–30 minutes into the simulations as the newly emerged and pre-existing ambient magnetic fields merge. The simulation A3 is initially very similar to A2, but only has a little flux emergence at later times, and the temperature remains of order 1.0 MK until late in the run when we see an increase toward 1.5 MK in the last 15 minutes.

2.2. Characteristics of Simulated TR and Coronal Lines

Let us now consider some examples of simulated spectral lines formed in the TR and corona of these models. We have chosen lines that have been previously well observed with space borne instruments such as Solar and Heliospheric Observatory (SOHO)/SUMER (Wilhelm et al. 1995) or are currently observable with Hinode/EIS (Culhane et al. 2007).
These lines cover a wide span of temperatures throughout the outer solar atmosphere and are summarized in Table 1.

The synthesized intensities $I_e$ in the various EUV spectral lines mentioned above are computed assuming that the lines are optically thin and the ionization state is in equilibrium

$$I_e = \frac{h\nu}{4\pi} \int \phi_e n_e n_H g(T_e) \, ds,$$

where the integration is carried out along the line of sight $s$, or in this case, vertically through the box along the z-axis using the variables as computed in the MHD code. Here, $n_e$ and $n_H$ represent the electron and hydrogen particle densities, respectively, $h\nu$ is the energy of the transition, and $g(T_e)$ is a function sharply peaked around the temperature of line formation defined by the atomic parameters describing the line excitation and ionization. In the following, we assume that the plasma and electron temperatures are equal ($T = T_e$). For the lines considered in this study, hydrogen is completely ionized and $n_e$ is simply proportional to the density. The line profile is computed assuming Doppler broadening

$$\phi_e = \frac{1}{\pi^{1/2}\Delta \nu_D} \exp\left[-\left(\frac{\Delta \nu - \nu_0}{\Delta \nu_D}\right)^2\right],$$

where $\Delta \nu = \nu - \nu_0$ is the distance (in frequency) from the rest frequency of the line, $\nu$ is the velocity, and $\mathbf{n}$ is the direction along the line of sight. The thermal broadening profile corresponds to a width

$$\Delta \nu_D = \frac{\nu_0}{c} \sqrt{\frac{2kT}{m_A}},$$

where $m_A$ is the mass of the radiating ion. We use the CHIANTI database (Dere et al. 1997; Young et al. 2003) to determine the atomic parameters of the spectral lines.

Let us first examine how the line profiles vary as a function of time at a given typical location picked more or less at random from the B1 model. The TR line profiles, exemplified by the C iv and Si vii lines which are representative of the lower and upper TR, respectively, are shown in Figure 3. These lines show a fairly similar evolution with significant variations in line intensity, shift, and width on timescales down to less than a minute. We see strong redshift events accompanied by brightening such as at 18 minutes and at 56 minutes, and a weaker blueshifted event at 47 minutes. It is clearly evident that the bulk of the C iv line is redshifted at almost all times, while the Si vii line has little obvious net shift in either direction.

The coronal line profiles, exemplified by Fe xii and Fe xv and also calculated from the B1 model, are shown in Figure 4. In addition for these lines, we find large variations in the line parameters on several timescales, including those shorter than 1 minute. The coronal lines become brighter as the corona grows hotter at times later than 20 minutes in the simulation. The strong brightening and blueshift seen at 47 minutes is found to be associated with a strong heating event (the Joule heating per particle in the B1 model at this time is shown in Figure 9, where the event causing the line variation is clearly visible as a tongue of heating stretching far into the corona). Note that the coronal lines at this location are preferentially blueshifted.

We can also examine the properties of the lines by examining the lower order moments: the total intensity, line shift, and line width. In Figure 5, we show the line moments of the typical lower TR C iv line and in Figure 7 the coronal Fe xii line, both at time $t = 2700$ s in the high field strength A2 run. At this time, the magnetic field in the corona forms two sets of loops connecting footpoints that are located in two bands centered roughly at $x = 7$ Mm and $x = 13$ Mm as shown in Figure 6. The field lines in the center of the figure have rotated compared with their original orientation parallel to the $x$-axis at the beginning of the simulation and are now oriented at roughly $45^\circ$ between the band of footpoints located near $x = 7$ Mm and the band located near $x = 13$ Mm (see also Martínez-Sykora et al. (2009) for a discussion of the interaction of pre-existing and newly emerged fields in the corona).

The left panels of Figures 5 and 7 show the total intensity, the middle panel the line shift, and the right panel the line width. There is no clear correlation between the C iv intensities seen in Figure 5 and the location of the footpoints seen in Figure 6. Rather, C iv emission is finely structured on spatial scales as small as a few hundred kilometers and as large as extended linear structures that are up to several megameters long. The contrast in the intensity is very high in this TR emission; the maximum intensity is 2 orders of magnitude greater than the average intensity. While the intensities do not obviously reflect the topology of the magnetic field, the line shifts and widths are more clearly related to the field. The Doppler shifts show intense redshifts and blueshifts concentrated to positions where the field comes up vertically from the chromosphere below, i.e., in the loop footpoints, and seem to form linear structures oriented at $45^\circ$ between the footpoints located at $x = 7$ Mm and $x = 13$ Mm. Similarly, there is a tendency for the greatest line widths to be found in the footpoint bands, while the line width generally appears narrower in the regions in between the footpoints. However, note that we also find some large line width regions stretching in thin linear structures between the footpoint bands and aligned with the field lines.

The pattern of intensities seen in the coronal lines is quite different from that of the TR emission, as seen for example in the Fe xii line intensities shown in Figure 7. In this line the coronal magnetic field, largely stretching between the footpoint bands, is clearly delineated as bright structures oriented at $45^\circ$ between the footpoints at $x = 7$ Mm and $x = 13$ Mm. The intensity contrast between maximum and average intensity is not as high as for the TR lines; it is of the order of 10 for the Fe xii line. The line shifts quite evidently show that blueshifts are concentrated to the footpoint bands with quite high amplitudes (the maximum blueshift is $-75$ km s$^{-1}$ while the maximum redshift is only $35$ km s$^{-1}$). Note that footpoint intensities are much lower than in the bodies of the loops between the footpoints, presumably as footpoint temperatures are too low to give much emission in coronal lines. The lines are both redshifted and blueshifted in the loop bodies in between the footpoints. The line widths are clearly greatest in bands along the loop footpoints, but again

| Line ID | $\lambda$ (nm) | log($T_{max}$) | Mass (amu) |
|---------|----------------|----------------|------------|
| He ii   | 25.63          | 4.9            | 4.0        |
| C iv    | 154.82         | 5.0            | 12.0       |
| O vi    | 103.19         | 5.5            | 16.0       |
| Si vii  | 27.54          | 5.8            | 28.1       |
| Fe xii  | 19.51          | 6.1            | 55.9       |
| Fe xiv  | 27.4           | 6.2            | 55.9       |
| Fe xv   | 28.42          | 6.3            | 55.9       |

Table 1 Simulated Lines
Figure 3. Simulated intensity profiles of the C\textsc{iv} and Si\textsc{vii} lines as functions of time from the low-resolution model B1. Example line profiles at selected times (at $t = 500, 1500, 2500, 3500$ s) are shown in the lower panels. The C\textsc{iv} line is formed in the lower TR at 10$^5$ K and has been extensively observed by the high-resolution telescope spectrometer and \textit{SOHO}/SUMER instruments. The Si\textsc{vii} line is formed in the upper TR at 6.3 × 10$^5$ K and has been observed by \textit{Hinode}/EIS.

Figure 4. Simulated intensity profiles of the Fe\textsc{xii} and Fe\textsc{xv} lines as functions of time from the low-resolution model B1. Example line profiles at selected times (at $t = 500, 1500, 2500, 3500$ s) are shown in the lower panels. The Fe\textsc{xii} line is typical for the low temperature corona and is formed at 1.25 MK. The Fe\textsc{xv} line is formed in hotter plasma at temperatures around 2 MK. Both lines are well observed with the \textit{Hinode}/EIS instrument.
Figure 5. Total intensity (left), Doppler shift (middle), and line width (right) in the $C_{\text{iv}}$ line at time $t = 2700$ s in the A2 simulation. The velocity scale is from $-40$ km s$^{-1}$ (blue) to 40 km s$^{-1}$ (red). Line widths range from narrow black to wide yellow/red with a maximum of 51 km s$^{-1}$. The average line shift in the $C_{\text{iv}}$ line is 6.6 km s$^{-1}$ (redshift is positive).

Figure 6. Vertical magnetic field $B_z$ (gray-scale image) in the photosphere and magnetic field lines that reach the corona (blue) at $t = 2700$ s in the A2 simulation.

we find some regions of large line width aligned with high intensities stretching along a subset of loops.

To quantify the coronal velocity pattern further, we plot normalized histograms for the Doppler shift and line width of the $Fe_{\text{xiv}}$ line, again from time $t = 2700$ s in simulation A2. In constructing the histograms, we averaged the line emission over a typical *Hinode* EIS resolution element of 1.5 Mm. It is interesting to compare the result with Figure 3 of Hara et al. (2008) who plot equivalent histograms constructed from observations of active region 10938 on 2007 January 18 for the same $Fe_{\text{xiv}}$ line. Figure 8 confirms that this line is strongly shifted to the blue, on average 10 km s$^{-1}$, in the footpoints of the simulation while showing less, if any, shift in the bulk of the emission. This fits very well with the findings of Hara et al. The line widths in the simulation are also on average greater in the loop footpoints, but both the absolute value of the widths and the difference between the footpoints and the loop bodies are smaller in the simulations than that found in the observations. On the Sun, there are presumably much larger motions on smaller scales (e.g., from turbulence) than those produced in the simulations.

3. CORONAL HEATING IN THIS MODEL

Let us now put the simulated “observed” line profiles described into context with the coronal heating that characterizes these models.

We find that in all the models considered the Joule heating per particle ($\propto j^2/\rho$) is concentrated to certain regions and structures that, in general, follow the magnetic field lines. This is illustrated in Figure 9 which shows the heating per particle in the B1 simulation at time 2810 s; a time of vigorous heating as emerging flux enters the corona and proceeds to reconnect with the pre-existing ambient field. This heating is highly irregular both in space and time, but is in large part concentrated near the upper chromosphere, TR, and lower corona, i.e., to a height range spanning from roughly $z = 1.5$ Mm to $z = 5$ Mm. Note that regions of high heating (colored in green–blue/purple) in part are aligned along loop like structures, in part in linear structures that stretch high into the corona. In addition, we find a weaker component (colored in red with high transparency) which, though it also follows the structure of the magnetic field, seems much more evenly spread out and space filling than the regions of (sporadic) high heating. In time, the linear regions of high heating are short lived, of order 100 s or shorter. The linear structures sometimes also move horizontally over a thousand kilometers or so during their lifetime. The loop-shaped structures are longer lived, though the heating varies continuously in time also here (see the animations provided in the online journal).

High heating events are found to occur in spatially localized bursts that lead to increases in the temperature but also large velocities in these models. The number and amplitude of bursts increase with increasing photospheric field strength and also with newly emerging field into the corona. For example, the large linear structure that reaches high into the corona, seen in the middle of the left panel of Figure 9 is the event that is visible in Figure 3 as a line shift of the TR lines at time $t = 47$ minutes and that is also evident in Figure 4 as a line shift and, especially for the $Fe_{\text{ xv}}$ line, by an increase in the intensity.

The concentration of heating toward the upper chromosphere, TR, and lower corona can be understood by considering the various scale heights involved. We find that Joule heating caused
by the braiding of the magnetic field decreases exponentially with height from the upper photosphere with a scale height that is closely related to the scale height of the magnetic energy density ($B^2/2\mu_0$). This is the same result as found earlier in the Gudiksen & Nordlund (2005) model (see Figure 7 of that paper), where a heating scale height of order 650 km was found in the chromosphere, rising to 2.5 Mm in the low corona and 5 Mm in the extended corona. Similar heating scale heights are found in Martínez-Sykora et al. (2008, 2009). These scale heights are much greater than the chromospheric pressure scale height of 200 km while, on the other hand, being much smaller than the 50 Mm coronal scale height. Thus, in models where chromospheric/coronal heating is determined by photospheric work on the magnetic field—with or without flux emergence—we find that the heating per unit mass (or equivalently per particle) necessarily is concentrated toward the TR and the low corona: low in the chromosphere the particle density is very high and therefore the heating rate per particle is low. At the same time, radiative losses are very efficient at these heights, thus the plasma easily rids itself of deposited energy via radiation. However, as long as the scale height of heating

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**Figure 7.** Total intensity (left), Doppler shift (middle), and line width (right) in the Fe xii line at time $t = 2700$ s in the A2 simulation. The velocity scale is from $-40$ km s$^{-1}$ (blue) to 40 km s$^{-1}$ (red). Line widths range from narrow black to wide yellow/red with a maximum of 76 km s$^{-1}$. The average line shift in the Fe xii line is $-2.7$ km s$^{-1}$ (blueshift is negative).

**Figure 8.** Normalized histograms for the Fe xiv 27.4 nm line from the A2 simulation at time $t = 2700$ s. Doppler velocity (left) and line width (right). Thin lines show histograms made from the entire area of the simulation and thick lines display regions where the field is nearly vertical at $z = 2$ Mm above the photosphere, i.e., loop footpoints. Dotted lines are the same as the thick lines with a different normalization.

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**Figure 9.** Current density squared per particle ($j^2/\rho$) at time $t = 2810$ s in simulation B1 (red weakest, blue/purple strongest, regions of very low heating are given high transparency and are therefore invisible) and vertical magnetic field $B_z$ in the photosphere (the range in field strength is set to saturate at ±1500 G), side view (left) and top view (right). The large linear structure that reaches high into the corona, seen in the middle of the left panel, is associated with strong line shifts at $t = 47$ minutes in Figure 4.

(Animations of this figure are available in the online journal.)
Figure 10. Average Joule heating $\eta j^2/\rho$ (solid) and average radiative losses $L_r = -n_e n_H f(T)/\rho$ (dot-dashed) per unit mass in the chromosphere and lower TR (upper panel) and corona (lower panel) in simulation B1. The average temperature is overplotted for reference (dashed).

Figure 11. Average line shift in the B1 model as a function of time for the lines of Table 1 (upper panel): He ii (dotted), C iv (solid), O vi (dashed), Si vii (dot-dashed), and Fe xii (triple-dot–dashed). The lower panel shows the average line shifts and the average velocity (dashed) as a function of log temperature in simulation B1. The error bars in the average line shift are given by the standard deviation in the average velocities as a function of time.

At some point this heating overwhelms the plasma’s ability to radiate, and the temperature rises to coronal temperatures, as illustrated in Figure 10. This figure shows the average Joule heating ($\eta j^2$) and radiative cooling ($L_r = -n_e n_H f(T)$) per unit mass as a function of height in the model B1. The temperature continues to rise until thermal conduction becomes effective enough to balance the deposited energy at temperatures approaching 1 MK. This temperature is high enough that the density scale height (~50 Mm) becomes much larger than the heating scale height and the heating per particle therefore decreases exponentially, roughly as the heating rate (or the magnetic energy density) itself.

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4. DISCUSSION AND CONCLUSIONS

Simulations of this sort lead to a coronal heating model where sporadic heating of material occurs mainly in the vicinity of the upper chromosphere, TR, and lower corona. When a heating event occurs in the upper chromosphere or TR, cooler material is rapidly heated to coronal temperatures. When heating occurs in the lower corona, the thermal conduction increases rapidly with rising coronal temperatures and the cooler chromospheric material just below can be heated by a thermal wave. It is our contention that these processes happen sufficiently rapidly that material is heated more or less “in place” before it has time to move. Thus, newly heated TR material will be characterized by higher than hydrostatic pressures, and we expect, on average, to find a plug of high pressure material that seeks to expand into regions of lower pressure on either side: downward toward the chromosphere and upward toward the corona.

The flows away from the high pressure plug result in the average velocities found in the TR and lower corona as shown in Figure 11. Average line shifts found in simulation B1 are relatively constant, showing a large apparent downflow in the TR and a smaller apparent upflow at coronal temperatures. These line shifts represent an actual material flow, on average, material is flowing away from the temperature of some $7 \times 10^5$ K with a maximum downward velocity of some $5 \text{ km s}^{-1}$ at $3 \times 10^4$ K and a maximum upward velocity of $2 \text{ km s}^{-1}$ at 1 MK as shown in the lower panel of Figure 11. The simulated line shifts reflect this flow trend fairly accurately, but due to the weighting of the line profile with density ($\propto n_e n_H$) the line shifts are exaggerated in the TR. We find that the average line shift in the coolest TR lines, He ii and C iv, are of order 7–8 km s$^{-1}$. The O vi line, formed near $3 \times 10^5$ K, is also shifted, but less, with an average that lies near 4 km s$^{-1}$. The upper TR Si vii line is initially blueshifted, but as time passes in the simulation the line shift
decreases toward zero and is slightly redshifted at the end of the run. Finally, the Fe xii line has an average line shift of zero initially but is slightly blueshifted at the end of the simulation run. These characteristics are also found in the higher resolution A2 and A3 models with high magnetic field strengths, which both show line shifts with amplitudes very similar to those found in the lower resolution B1 model. It is only the relatively weak field A1 model that differs markedly in terms of average line shifts as discussed below.

Note that although a net redshift is observed in TR lines which corresponds to a material (downward) flow out of the TR, the coronal mass is actually constant or growing in these simulations. Thus, material does not flow into the corona through the TR. Rather, material is heated, in place and rapidly, to higher temperatures. This heated material leads to an overpressure compared to hydrostatic equilibrium, and excess material is subsequently pushed out of the upper TR upward into the corona and down toward the chromosphere.

The models presented here have horizontal spatial resolution of 65 km and 32.5 km and vertical resolution of 32 km and 28 km for the “low-” and “high-” resolution models, respectively. One could reasonably ask how well the simulated line profiles reflect the underlying dynamics of the atmosphere and furthermore how well the dynamics themselves are modeled, given the relatively small number of computational points that cover the TR. We find that TR dynamics, in particular those underlying the persistent redshifts, cover scales that are larger than those defining the line forming regions and conclude that the simulated line shifts and intensities are faithfully reproducing these dynamics; i.e., average line shifts are due an average plasma flow as shown in Figure 11. On the other hand, we do believe that as the spatial resolution of these models becomes better the structure of the velocity fields will become finer, with increasing line widths as measured at a given spatial scale as a result. The effect of poor spatial resolution on the structure and dynamics of the atmosphere itself is difficult to ascertain and the ultimate test of the importance of spatial resolution on the TR will have to await the completion of very large 3D models. However, we have compared 1D models using the BIFROST code with resolutions of 28, 14, and 7.5 km in the TR without finding large quantitative nor qualitative differences in the temperature structure or in the conductive heat flux. This also remains true on comparison with 1D models with adaptive grids (e.g., Hansteen 1993) where the spatial resolution is much better, of the order of 125 m at the sites of the largest temperature gradient.

The arguments above are strengthened by considering the A1 model which has a vertical spatial resolution of 28 km, but where we do not find redshifted TR spectral lines. In this model, which has significantly less coronal heating than that found in the other three models, we do not find average flows out of the upper TR. Figure 12 shows the average flow velocity for plasma at \( T = 10^5 \) K and \( T = 3 \times 10^5 \) K; the average varies as a function of time as acoustic waves pass into the corona, but the time average flow is very small; we find 1.2, 0.8, −2.0, and −3.1 km s\(^{-1}\) for temperatures of \( T = 30, 100, 300, \) and 500 K, respectively. This weak field, low heating model thus supports our scenario that, in the models presented, the TR redshifts are a direct consequence of vigorous heating to coronal temperatures at low heights, with the redshifts disappearing as the amplitude of the heating decreases.

Further support for the connection between a high pressure TR and redshifts is shown in Figure 13 where we show the correlation between the redshift measured in the C iv TR line and the maximum downward oriented pressure gradient −\( dp/dz/\rho \) in the B1 model. The figure shows that high redshift occurs in regions with a large downward directed pressure gradient. Note that Figure 13 also shows that a fairly large population of large downward directed pressure gradients co-exist with line shifts that are close to zero. After comparing with the magnetic field vector at the location of these points, it is found that this population is representative of points where the magnetic field is mainly horizontal; i.e., where there is a significant contribution from the Lorentz force, presumably located in the body of low-lying cooler loops. This latter population does not have much impact on the average line shifts.

The proposed mechanism implies the ubiquitous presence of short-lived (∼100 s) heating events during which high speed flows for a wide range of temperatures are expected to occur. It is tempting to speculate that these flows in the models may be related to the high speed upflows that have been deduced from significant blueward asymmetries in Hinode/EIS observations of many TR and coronal lines (Hara et al. 2008; De Pontieu et al. 2009; McIntosh & De Pontieu 2009a, 2009b). However, the current models still have fairly low spatial resolution (which may lead to underestimates of the amplitude of individual high
speed flows since the released energy necessarily is spread out over larger volumes), so a detailed comparison with such observations will be the subject of a future paper.

The sporadic heating of the TR material as outlined here also has consequences for the intensities of lines formed in the chromosphere–corona interface, and thus for the shape of the DEM curve. The rise in the DEM toward lower TR temperatures indicates that much more plasma is emitting in this temperature range than is accounted for by models where the TR structure is dominated by thermal conduction from overlying coronal structures. A straightforward extension of this work in which simulated DEMs are calculated is planned, where we will also compare intensities directly with those observed with SOHO/SUMER and Hinode/EIS.

In summary, we find that outflows from the upper TR can be explained as a natural consequence of episodically heated models where the heating per particle is concentrated toward the lower atmosphere, from the upper chromosphere to low corona. Chromospheric material is heated in place to coronal temperatures, or rapidly heated coronal material just above the TR leads to a thermal conduction front heating the chromospheric material just below the Joule dissipation region to coronal temperatures; both variants lead to a high pressure plug of material at upper TR temperatures that relaxes toward equilibrium by expelling material. We have further shown that heating with these characteristics naturally arises when coronal heating is assumed to be caused by the dissipation resulting from the braiding of the photospheric magnetic field, potentially aided by the emergence of new magnetic flux into the corona. This mechanism leads to Joule heating that has an exponential decrease with height closely related to the scale height in the magnetic energy density, a scale height that is longer than the chromospheric density scale height, but shorter than the corresponding coronal density scale height, forcing the heating per particle to be greatest in the vicinity of the TR.

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