NUMERICAL SIMULATION OF SOLAR MICROFLARES IN A CANOPY-TYPE MAGNETIC CONFIGURATION

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ABSTRACT

Microflares are small activities in the solar low atmosphere; some are in the low corona while others are in the chromosphere. Observations show that some of the microflares are triggered by magnetic reconnection between the emerging flux and a pre-existing background magnetic field. We perform 2.5-dimensional, compressible, resistive magnetohydrodynamic simulations of the magnetic reconnection with gravity considered. The background magnetic field is a canopy-type configuration that is rooted at the boundary of the solar supergranule. By changing the bottom boundary conditions in the simulation, a new magnetic flux emerges at the center of the supergranule and reconnects with the canopy-type magnetic field. We successfully simulate the coronal and chromospheric microflares whose current sheets are located at the corona and the chromosphere, respectively. The microflare with a coronal origin has a larger size and a higher temperature enhancement than the microflare with a chromospheric origin. In the microflares with coronal origins, we also found a hot jet (\(1.8 \times 10^6\) K), which is probably related to the observational X-ray/\(\alpha Ca\) bright point in the microflares that have chromospheric origins. The study of parameter dependence shows that the size and strength of the emerging magnetic flux are the key parameters that determine the height of the reconnection location, and they further determine the different observational features of the microflares.

Key words: magnetic reconnection – magnetohydrodynamics (MHD) – methods: numerical – Sun: flares

Online-only material: color figures

1. INTRODUCTION

Microflares, or subflares, which are small-scale and short-lived solar activities, have been studied since the past century (Smith & Smith 1963; Svestka 1976; Tandberg-Hanssen & Emslie 1988). Their typical size, duration, and total released energy are 5–20 arcsec, 10–30 minutes, and \(10^{26}–10^{29}\) erg (Shimizu et al. 2002; Fang et al. 2006, 2010), respectively. Microflares have been observed in many wavelengths including \(H\alpha\) (Schmieder et al. 1997; Chae et al. 1999; Tang et al. 2000), extreme ultraviolet (EUV; Emslie & Noyes 1978; Porter et al. 1984; Chae et al. 1999; Brosius & Holman 2009; Chen & Ding 2010), soft X-ray (SXR; Golub et al. 1974, 1977; Tang et al. 2000; Shimizu et al. 2002; Kano et al. 2010), hard X-ray (HXR; Lin et al. 1984; Qiu et al. 2004; Ning 2008; Brosius & Holman 2009), and microwave (Gary & Zirin 1988; Gopalswamy et al. 1994; Gary et al. 1997). However, not all microflares have emissions at all wavelengths. Observations show that most of the bright X-ray microflares also appear at EUV and \(H\alpha\) bands, but only part of the \(H\alpha\) microflares have their counterparts in X-ray emission (Zhang et al. 2012).

Observational characteristics of microflares, such as heating, relation with the magnetic field, duration, and coincidence between the different wavelengths, imply that microflares are produced by a magnetic reconnection process that is similar to big flares. For example, Qiu et al. (2004) found that about 40% of microflares show HXR emissions at more than 10 keV and they show microwave emissions at about 10 GHz, all of which are typical features of flares. Recently, Ning (2008) found that roughly half of microflares display the Neupert effect as previously revealed in flares. On the other hand, the other microflares may have their origins in the lower atmosphere. Brosius & Holman (2009) found that microflares are bright in the chromospheric and transition region spectral lines, which is consistent with chromospheric heating by nonthermal electron beams. Jess et al. (2010) concluded that the microflares in their study are due to magnetic reconnection at a height of 200 km above the solar surface.

Because parts of microflares are located at emerging flux regions, they are probably due to magnetic reconnection driven by the new emerging magnetic flux (EMF) with a pre-existing magnetic field. Schmieder et al. (1997) found that the X-ray loops of microflares appear at the locations of the EMF. Chae et al. (1999) found some repeatedly occurring EUV jets where the pre-existing magnetic field was “canceled” by the new EMF with opposite polarity. Tang et al. (2000) indicated that the new EMF successfully emerged about 20 minutes before the peaks of the \(H\alpha\) and SXR brightenings. Half of the events studied by Shimizu et al. (2002) show the small-scale emergences of the magnetic flux loops in the vicinity of the transient brightenings. Kano et al. (2010) found that EMFs and moving magnetic features (MMFs) are related to the energy release for at least half of the microflares around a well-developed sunspot. Therefore, to understand the dynamics of the microflares, it is crucial to simulate the magnetic reconnection associated with an EMF.

The numerical simulations of the flux emergence scenario started in the late 1980s (Horiuchi et al. 1988; Matsumoto et al. 1988; Shibata et al. 1989a, 1989b). Later, more realistic and complicated numerical experiments (Magara 2001; Fan 2001; Manchester et al. 2004; Isobe et al. 2005; Leake & Arber 2006; Isobe et al. 2007; Hood et al. 2009) were presented to study the emergence of the twisted flux from the convection region to the solar corona. The simulations are mainly based on the instability proposed by Parker (1966), and the results of these simulations can account for the formation of a filament or newly active region. To study the solar eruptions, additional
two- and three-dimensional numerical experiments are focused on the interaction between the emerging flux and the pre-existing coronal magnetic configuration. Through the use of two-dimensional simulations, the magnetic reconnection due to the flux emergence can explain the solar jets or the explosive events (Yokoyama & Shibata 1995; Jin et al. 1996; Nishizuka et al. 2008; Ding et al. 2010, 2011); they can even trigger the onset of coronal mass ejections (Chen & Shibata 2000). Further three-dimensional simulations gave the same but more detailed results, revealing that the emerging flux plays a key role in the formation of many solar activities, such as type II spicules, erupting filaments or flux ropes, Ellerman bombs, and so on (Galsgaard et al. 2005, 2007; Török et al. 2009; Archontis Török 2008; Archontis & Hood 2009; Martínez-Sykora et al. 2011).

From the description of previous works, it is seen that the EMF is responsible for many different phenomena, and one main cause of the diversity of the dynamics is that the reconnection occurs at different heights in the solar atmosphere (Shibata 1996; Chen et al. 1999). Our simulation in this paper is mainly focused on the microflares in a canopy-type magnetic configuration, where the height of the reconnection is self-consistently determined by the EMF. In our previous papers (Jiang et al. 2010; Xu et al. 2011), we presented a 2.5-dimensional magnetohydrodynamic (MHD) simulation of the magnetic reconnection in the chromosphere that qualitatively reproduces the temperature enhancement observed in the chromospheric microflares. However, the magnetic configuration in that simulation, which is identical to the one in Chen et al. (2001), is too idealized and the reconnection site was specified at the computational center. Moreover, the corona was not included, so the simulations cannot provide any information about the EUV and SXR features. In this paper, we adopt a more realistic magnetic configuration and extend the solar atmosphere up to the corona to see what determines whether a microflare will be of coronal origin or chromospheric origin and compare the difference between the two types of microflares. The paper is organized as follows: the numerical method is described in Section 2, the numerical results are presented in Section 3, and the discussion and summary of the results are given in Section 4.

2. NUMERICAL METHOD

2.1. Basic Equations

Keeping the magnetic field divergence-free is one of the biggest problems in all of the MHD codes. Because of discretization and numerical errors, the performance of the MHD code may be unphysical (Brackbill & Barnes 1980). There are several ways to maintain ∇ · B = 0 for MHD equations: (1) the 8-wave formulation (Powell et al. 1999), (2) the constrained transport method (Evans & Hawley 1988; Stone & Norman 1992), and (3) the projection scheme (Brackbill & Barnes 1980). The comparison between these methods showed that different methods have their own advantages and disadvantages (Tóth 2000). Moreover, one can rewrite the original MHD equations using the vector potential A instead of the magnetic field B or using the vector magnetic potential or the Euler potential. The advantage is that the divergence-free condition is always satisfied; however, the MHD equation should be rewritten. Our method, as introduced in Jiang et al. (2012), adopts the 8-wave method. As suggested by Dedner et al. (2002), we use the extended generalized Lagrange multiplier (EGLM) MHD equations rather than pure MHD equations, which include two additional waves to transfer the numerical error of ∇ · B. The local divergence error can be damped and passed out of the computational domain. The adopted dimensionless EGLM-MHD equations that include resistivity and gravity are given as follows:

\[ \frac{\partial \rho}{\partial t} + \nabla \cdot (\rho \mathbf{v}) = 0, \quad (1) \]

\[ \frac{\partial (\rho \mathbf{v})}{\partial t} + \nabla \cdot \left( \rho \mathbf{v} \mathbf{v} - \mathbf{B} \right) = - \nabla \cdot (\mathbf{B} + \rho \mathbf{g}), \quad (2) \]

\[ \frac{\partial \mathbf{B}}{\partial t} + \nabla \times (\mathbf{vB} - \mathbf{Bv} + \psi \mathbf{I}) = - \nabla \times (\eta \nabla \times \mathbf{B}), \quad (3) \]

\[ \frac{\partial e}{\partial t} + \nabla \cdot \left( \mathbf{v} \left( c^2 + \frac{1}{2} B^2 - \frac{\rho}{\gamma - 1} \right) - \mathbf{B} \cdot \mathbf{v} \right) = - \mathbf{B} \cdot (\nabla \psi) - \nabla \cdot \left( (\eta \nabla \times \mathbf{B}) \times \mathbf{B} \right) + \rho \mathbf{g} \cdot \mathbf{v}, \quad (4) \]

\[ \frac{\partial \psi}{\partial t} + c^2 \frac{\nabla}{c_p^2} \cdot \mathbf{B} = - \frac{c_p^2}{c_p^2} \psi, \quad (5) \]

where eight independent conserved variables are the density (\( \rho \)), momentum (\( \rho \mathbf{v}, \rho \mathbf{v}, \rho \mathbf{v} \)), magnetic field (\( B_x, B_y, B_z \)), and total energy density (\( e \)). The expression of the total energy density is \( e = \rho / (\gamma - 1) + \rho \mathbf{v}^2 / 2 + B^2 / 2 \), where \( \gamma = 1.1 \) is taken in all of our computations. The pressure \( p \) and the temperature \( T \) are dependent on the eight conserved variables, \( \mathbf{g} \) is the gravity vector, and \( \eta \) is the magnetic resistivity coefficient. Finally, \( \mathbf{I} \) is the unity matrix. The main variables are normalized by the quantities given in Table 1.

In the EGLM-MHD equations, \( \psi \) is a scalar potential propagating the divergence error, \( c_h \) is the wave speed, and \( c_p \) is the damping rate of the wave (Dedner et al. 2002; Matsumoto 2007). As suggested by Dedner et al. (2002), the expressions for \( c_h \) and \( c_p \) are

\[ c_h = \frac{c_{dA}}{\Delta t} \min(\Delta x, \Delta y), \quad (6) \]

\[ c_p = \sqrt{-\left( \delta t \frac{c_{dA}^2}{\ln c_{dA}} \right)}, \quad (7) \]

where \( \Delta t \) is the time step, \( \Delta x \) and \( \Delta y \) are the grid sizes, and \( c_{dA} \) is a safety coefficient less than 1. \( c_{dA} \in (0, 1) \) is a problem-dependent coefficient used to decide the damping rate for the waves of the divergence errors. We can see that \( c_h \) and \( c_p \) are not independent of the grid resolution and the scheme used. Hence, we must adjust their values for different situations.
2.2. Initial Condition

The computational box is located in the $x$–$y$ Cartesian plane. The $x$-axis is parallel to the solar surface, while the $y$-axis is perpendicular to the photosphere. As shown in Figure 1, the computational domain is $-100 \leq x \leq 100$ and $0 \leq y \leq 200$, where the length unit is $L_0 = 301.6$ km. In the simplified chromosphere and photosphere, the temperature is set to be uniform with the value of 6.6 after being normalized by $T_0 = 10,000$ K. This layer extends from the bottom to $y = 8.5$. After a thin transition region with a thickness of 1, the plasma temperature rapidly rises to 100 in the corona, which corresponds to 1 MK. Given the temperature distribution, we can obtain the density and pressure distributions according to the hydrostatic equilibrium. The initial distributions of the density, gas pressure, plasma beta ($\beta$), and temperature ($T$) are outlined by Figure 2.

As shown in Figure 1, two canopy-shaped magnetic structures are separated by a distance of 100 (corresponding to 30 Mm) to mimic the two boundaries of a supergranular cell. The canopy magnetic field, which was originally proposed by Gabriel (1976) and Giovanelli (1980, 1982) for studying the chromospheric magnetic field, is rooted at the supergranular boundaries. To create a canopy magnetic field, we introduce several “magnetic charges” located beneath the photosphere. According to Gauss’s law for magnetism, the magnetic field (at the position $\mathbf{r}$) generated by the “magnetic charges” (at the position $\mathbf{r}'$) can be written as $\mathbf{B} = c_B (\mathbf{r} - \mathbf{r}') / |\mathbf{r} - \mathbf{r}'|^2$ in two-dimensional geometry ($c_B$ is a constant related to the strength of the field) and $\mathbf{B} = c_B (\mathbf{r} - \mathbf{r}') / |\mathbf{r} - \mathbf{r}'|^3$ in three-dimensional geometry, which is similar to the electric field. The locations of the “magnetic charges” are depicted in Figure 1. The magnetic configuration in our simulation is not exactly the same as the canopy model proposed by Gabriel (1976) and Giovanelli (1980, 1982).

2.3. Resistivity Model

To reproduce fast reconnection (Petschek 1964), we adopt a nonuniform anomalous resistivity. This kind of resistivity may be due to some microscopic instabilities; however, it is still not clear how these instabilities drive the macroscopic reconnection. The mathematical form of the assumed anomalous resistivity model (Ugai 1985; Chen & Shibata 2000; Yokoyama & Shibata 2001) is

$$\eta = \min \left( \left( \frac{|j|}{j_c} - 1 \right), \eta_{\text{max}} \right) \text{ for } |j| \geq j_c, \quad (8)$$

where $|j|$ is the total current density, $j_c$ is the threshold above which the anomalous resistivity is excited, and $\eta_{\text{max}}$ is the maximum value of the resistivity. As shown by Equation (8), the resistivity ($\eta$) is dependent on the current density. The anomalous resistivity ($\eta$) is switched off when $|j| < j_c$.

2.4. Boundary Condition

The magnetic flux emergence is realized by changing conditions at the lower boundary (Forbes & Priest 1984; Chen & Shibata 2000; Ding et al. 2010). In our case, $B_x$ and $B_y$, rather than the magnetic flux function in previous works, are independent variables; we directly change the value of the magnetic field at the boundary. The magnetic field ($B_x$, $B_y$, $B_z$) changes with time, while the other variables, e.g., $\rho, v, p$, are fixed to the initial values at this boundary. The mathematical form of the EMF is given by

$$B_x(x, y) = -B_e \frac{y - y_e}{l_e x^2 + y^2}, \quad (9)$$

$$B_y(x, y) = B_e \frac{x - x_e}{l_e x^2 + y^2}, \quad (10)$$

![Figure 1](image-url)  
**Figure 1.** Initial conditions for the magnetic configuration and the temperature distribution. The ranges of the computational box are from $-100$ to $100$ in the $x$-direction and from $0$ to $200$ in the $y$-direction. The shadowy region is the damping zone. The “magnetic charges” are located beneath the solar surface. Note that the “magnetic charges” are samples to show the positions of these “charges.” The units of length, temperature, time, and velocity are $301.6$ km, $10,000$ K, $33.1$ s, and $9.09$ km s$^{-1}$, respectively. (A color version of this figure is available in the online journal.)

![Figure 2](image-url)  
**Figure 2.** Initial distributions of the density ($\rho$), gas pressure ($p$), plasma beta ($\beta$), and temperature ($T$). The density, gas pressure, and plasma beta distributions use the left $y$-axis, and the right blue $y$-axis indicates the temperature. The units of density, gas pressure, length, temperature, time, and velocity are $2.56 \times 10^{-8}$ kg m$^{-3}$, $2.12$ N m$^{-2}$, $301.6$ km, $10,000$ K, $33.1$ s, and $9.09$ km s$^{-1}$, respectively. (A color version of this figure is available in the online journal.)
in the range $x^2 + y^2 < r_0^2$, $y < 0$, where $B_c$ is the strength of the EMF, $t_e$ is the end time of the emergence, $r_0$ is the half-width of the EMF, and $x_c$ and $y_c$ are the center coordinates of the EMF in the $x$- and $y$-directions, respectively. In our simulation cases, $x_c$ and $y_c$ are set to be 0 and $-3.6$, respectively. Equations (9) and (10) are applied to the magnetic field $B$ until $t = t_e$. Later, all of the physical values are fixed.

Free boundaries (equivalent extrapolation) are adopted at the upper side (i.e., $y = 200$). However, because of the low plasma beta ($\beta = 2 \rho / B^2$), the upper boundary becomes unstable when the jets or the waves propagate to this boundary. Thus, we add a damping zone from $y = 180$ to $y = 200$ in the $y$-direction as shown in Figure 1. Either the waves or the jets will be damped to the initial value as the formulae given below:

$$u(x, y) = u(x, y)D(y) + u_0(x, y)(1 - D(y)), \quad (11)$$

$$D(y) = \frac{1}{2} \left(1 - \tanh \left(\frac{6}{x_c - y_c} \left(y - y_c + \frac{y_c}{2} \right)\right)\right), \quad (12)$$

where $u$ represents variables $\rho$, $v$, and $p$ (not $B$); $u_0$ is the initial value; $D$ is the damping function; $y_c$ is the start position for the damping zone in the $y$-direction; and $y_e$ is the end position for the damping zone. In all of our cases, $y_c = 180$ and $y_e = 200$. Nonetheless, the damping zone may also lead to some reflections of the waves or the jets, but these reflections have little effect on the results since they are far away from the lower atmosphere that interests us. We use the fixed boundary condition for the left and right boundaries at $x = -100$ and $x = 100$, respectively.

2.5. Numerical Scheme and AMR Grid

The code used in our simulation is “MAP” (Jiang et al. 2012), which was developed by the solar group of Nanjing University. The MAP code is an Fortran code for the MHD calculation that uses adaptive mesh refinement (AMR; Berger & Oliger 1984; Berger & Colella 1989) and the Message Passing Interface parallelization. The MAP code also has three optional numerical schemes for the MHD part, namely, the modified MacCormack Scheme (Yu & Liu 2000), the Lax-Friedrichs scheme (Toth & Odstrcil 1996), and the weighted essentially nonoscillatory (WENO; Jiang & Wu 1999) scheme. In this paper, we mainly use the WENO scheme, which has a higher accuracy than the other two schemes. Moreover, it can keep the total variation diminishing (Harten 1997) property for the MHD part without any additional artificial viscosity. The base resolution of our simulation is $256 \times 256$ and the mesh refinement level is 4 in all of our cases. Thus, the effective resolution is $2048 \times 2048$ globally and the minimum grid size is 0.1 (about 30 km).

3. RESULTS

Our simulations show two typical cases: one with the reconnection occurring in the corona and the second occurring mainly in the chromosphere and partly in the transition region. The dynamics of the magnetic reconnection processes are similar in the early stages of both cases. Both show the following evolution: (1) an EMF emerges from the center of the supergranule; (2) the EMF reconnects with the pre-existing magnetic field; and (3) the temperature is enhanced near the X-point, and the inflow and outflow reconnections appear. However, the final results are very different. The differences between these two cases are presented in Sections 3.1 and 3.2.

3.1. Coronal Microflare

In this case, we set the parameters of the EMF to be $B_c = 32$ (corresponding to 520 G) and $\eta_{\text{max}} = 0.1$ and $r_0 = 8$ (corresponding to $\sim 2400$ km). The reconnection occurs when the local current density is larger than the threshold ($j_0 = 10$) in Equation (8). Figure 3 depicts the results of the simulation at different times. In this figure, the color stands for the temperature, solid lines represent the magnetic field, and arrows represent the velocities. The upper panel of Figure 3 shows that the EMF already rose into the photosphere and chromosphere. At time 55, the transition region has been pushed up to the height of $y = 35$ (i.e., about 10,000 km), and the fast reconnection has not yet happened. Later, when the magnetic reconnection rate reaches the maximum (time = 70, as shown by the upper left panel of Figure 4), a large amount of magnetic energy is released as the internal and kinetic energies. In this case, the reconnection occurs at the corona, and the size of this microflare is about 50 km $x = -40$ to 10 ($\sim 15,000$ km, i.e., $\sim 20$ arcsec). The hot ($\sim 1.8 \times 10^6$ K) and cold jets ($\sim 10^4$ K) formed around this time are similar to the results simulated by Yokoyama & Shibata (1995) and Nishizuka et al. (2008).

The reconnection rate and the one-dimensional distributions of the density, temperature, and $y$-component of the velocity along the red line in Figure 3 are all illustrated in Figure 4. The reconnection rate is calculated by the ratio of the inflow ($V_{\text{in}}$) and local Alfvén speeds ($V_A$). Both of the speeds are measured around the current sheet. The reconnection rate reaches the maximum around time $= 70$ (38 minutes), and its value is about 0.09, which is within the value of 0.01–0.1 predicted by Petschek (1964). The high reconnection rate partly results from some small plasmoid ejections from the X-point, which may increase the reconnection rate (Shibata et al. 1995; Shibata 1996). After time $= 76$, some numerical errors occur at the center of the EMF. Thus, the result after that time is unreliable, although the reconnection rate again increases (see the upper left panel of Figure 4). A roughly estimated lifetime of this coronal microflare is about 12 minutes if we consider that the duration is from the beginning (time $= 55$) to the end (time $= 75$) of the main reconnection. This is consistent with the observational values (10–30 minutes from others; Shimizu et al. 2002; Fang et al. 2006, 2010). The other three panels show the one-dimensional distributions of density, temperature, and $y$-component velocity at time $= 45$, respectively. As we mentioned above, a hot jet and a cold surge are formed in this case. The hot jet ($\sim 1.8 \times 10^6$ K), which originates from the reconnection region, is ejected into the higher corona with a velocity of 16 (about 140 km s$^{-1}$), corresponding to the observational EUV/SXR jet (Chae et al. 1999; Brosius & Holman 2009). The denser cold surge ($\sim 10^4$ K; the density is two orders of magnitude larger than the hot jet) is drawn to the corona by the hot jet. The cold surge later falls to the lower atmosphere with a speed of 3 ($\sim 30$ km s$^{-1}$). This process is similar to the observational Hα/Ca surge seen in Chae et al. 1999. The estimated size of the EUV/SXR or Hα bright point in our simulation is about 50 km $x = -40$ to 10 ($\sim 15,000$ km, i.e., $\sim 20$ arcsec).

3.2. Chromospheric Microflare

In this case, the strength and half-width of the EMF are $B_c = 12$ (200 G) and $\eta_{\text{max}} = 0.1$ and $r_0 = 6$ ($\sim 1800$ km). Compared with the coronal case, the reconnection process is similar, however, but the simulation result is totally different, since the canopy-type magnetic field prevents the weak EMF...
from emerging to a height as high as the coronal one. Therefore, the magnetic field accumulates in the chromosphere until the local current density exceeds the threshold ($\jmath_c = 10$, the same as the coronal case). Finally, we obtained a chromospheric microflare resulting from a fast reconnection mainly in the chromosphere and partly in the transition region. The results of this chromospheric microflare are shown in the three snapshots in Figure 5. Similar to Figure 3, the three panels correspond to the times before the reconnection, the time when the reconnection rate is increasing, and the time when the reconnection rate reaches the maximum (as shown by the upper left panel of Figure 6), respectively. In this case, the reconnection occurs at
Figure 4. Upper left panel shows the magnetic reconnection rate using the ratio between the inflow and Alfvén speeds. The three vertical dashed lines in this panel outline the times used in Figure 3. The other three panels are one-dimensional distributions of the density (RO), temperature (TE), and y-component velocity (VY) along the solid red line shown in the lower panel of Figure 3. The horizontal dashed lines show the initial values, and the vertical lines indicate the boundaries between the hot and cold jets in these three panels. The density, length, temperature, time, and velocity units are $2.56 \times 10^{-8}$ kg m$^{-3}$, 301.6 km, 10,000 K, 33.1 s, and 9.09 km s$^{-1}$, respectively.

the chromosphere, and the size of this microflare is about 15 from $x = -10$ to 5 ($\sim 4500$ km, i.e., $\sim 6$ arcsec). There are no obvious reconnection jets ejected into the corona, and we only obtain some temperature enhancements in the chromosphere and only partly in the transition region.

The reconnection rate and the one-dimensional distribution of the temperature along the red line in Figure 5 are both illustrated in Figure 6. The physical process is much simpler than what is seen in the corona case. As the reconnection occurs in the dense chromosphere, it is very difficult to heat the plasma to a very high temperature. The right panel of Figure 6 shows the temperature distribution at a height of 1500 km. The initial value of the temperature is 0.64 (6400 K), and the temperature enhancement is about 800 K at this layer. The low temperature region is due to the expansion of the EMF. As we know, the plasma satisfies the frozen-in condition in the solar atmosphere without the resistivity. Therefore, the plasma of the lower atmosphere is dominated by the strong magnetic field and finally such an adiabatic expansion will reduce the temperature. The low adiabatic index ($\gamma = 1.1$ in our simulations) can decrease such an effect, which also makes the simulated system similar to an isothermal process. Although we can see two regions where the temperature increased, as seen in Figure 6, these regions are not observed as two Hβ bright points; this is because the hot region is a shell covering the EMF, as seen from the lower panel of Figure 5. That is to say, we can only observe one Hβ bright point if the line of sight is along the y-axis. The size of this bright point is about 15 from $x = -10$ to 5 ($\sim 4500$ km, i.e., $\sim 6$ arcsec). There is no obvious temperature enhancement in the corona. Thus, this case can only explain the microflares with an Hα emission. The inflow (downward) speed in the lower panel of Figure 5 is about 6 km s$^{-1}$, which may be observed as a redshifted Doppler velocity in the coronal lines. The outflow velocity of the reconnection flow in the chromosphere is about 20 km s$^{-1}$, but the direction of this outflow is redirected by a background, canopy-type field. Eventually, this coronal outflow ($\sim 4$ km s$^{-1}$) moves upward next to the downward inflow as shown by the lower panel of Figure 5.

3.3. Parameter Dependence

So far, we have simulated one coronal case and one chromospheric case. An important question to answer is which parameters determine whether reconnection happens in the corona or in the chromosphere when an EMF appears. In this subsection, we study three parameters, i.e., the strength of the EMF ($B_e$), the maximum value of resistivity ($\eta_{\text{max}}$), and the half-width of the EMF ($r_0$). To understand the effects of the different parameters,
Figure 5. Temperature distributions (color scale), projected magnetic field (solid lines), and velocity field (vector arrows) at times 20 (upper panel), 40 (middle panel), and 72 (lower panel). The units of length, temperature, time, and velocity are 301.6 km, 10,000 K, 33.1 s, and 9.09 km s\(^{-1}\), respectively.

(A color version of this figure is available in the online journal.)

we perform extensive simulations by varying one parameter with the others remaining fixed.

Figure 7 shows the dependence of the reconnection process on the different strengths of the EMF magnetic field. Four values of \(B_e\) are studied, i.e., \(B_e = 12, 22, 32,\) and 42, whereas the other two parameters are fixed, i.e., \(\eta_{\text{max}} = 0.1\) and \(r_0 = 8\).

With the increasing of the strength of the EMF, the magnetic reconnection rate can reach the maximum much faster, and the maximum reconnection rate becomes increasingly larger when we set a stronger initial magnetic field. The right panel of Figure 7 shows how the maximum temperature and the height of the current sheet center change with the strength of the EMF.
magnetic field. The Max Temperature and the Height are very sensitive to the parameter $B_e$. A stronger EMF can result in a higher altitude that can generate a stronger current density. Thus, it is easy to trigger the reconnection process at an earlier time. The reconnection process is faster and the outflow is hotter than a weak EMF. When we set $B_e = 12$, we find that the height is 12.5 (3750 km), which is located nearly above the transition region. If we keep reducing the value of $B_e$, the reconnection will occur in the chromosphere.

Figure 8 shows the dependence of the various quantities on the different values of maximum resistivity ($\eta_{\text{max}}$). Four different values are selected, i.e., $\eta_{\text{max}} = 0.01, 0.05, 0.1, \text{and} 1.0$. The other two parameters are fixed, i.e., $B_e = 32$ and $r_0 = 8$. We achieved similar results to those described in our previous paper (Jiang et al. 2010). The variation of the resistivity value does not change the results too much. The reconnections occur almost around the height of 28 (8500 km) and the temperature enhancement occurs at a height of about 350 ($3.5 \times 10^6$).

The dependence on the different half-widths ($r_0 = 4, 6, 8,$ and 12) is depicted in Figure 9. The other two fixed parameters are $B_e = 32$ and $\eta_{\text{max}} = 0.1$. From the reconnection start time to the maximum time, a bigger EMF can rise to a higher altitude. We know that the plasma $\beta$ is lower in the higher altitude; thus, the magnetic reconnection can be very fast with the same resistivity. At the same time, the strength of the magnetic field of the EMF becomes weaker and weaker after expanding to such a long distance, so the temperature enhancement becomes smaller and smaller, as indicated by Figure 7. Therefore, the left panel of Figure 9 shows the reconnection rate; the reconnection duration increases with the increasing half-width of the EMF. In the right panel, we find that the height of the X-point is also sensitive to the size of the EMF. Combining the small magnetic strength of the EMF, we can get the result of the chromospheric microflare described in Section 3.2.

4. DISCUSSION AND SUMMARY

In this paper, we simulate the microflares produced by EMFs. The EMFs with different strengths and sizes lead to different results. The two types of microflares in our simulations, i.e.,
Figure 8. Left panel shows the dependence of the $\frac{\text{Max} V_{\text{in}}}{V_A}$ and the Max Rate Time on different maximum values of resistivity ($\eta_{\text{max}}$). The right panel depicts the dependence of the Max Temperature and the Height on different maximum values of resistivity ($\eta_{\text{max}}$). The length, temperature, time, and resistivity units are $301.6 \text{ km}$, $10,000 \text{ K}$, $33.1 \text{ s}$, and $3445.1 \text{ \Omega m}$, respectively.

Figure 9. Left panel shows the dependence of the $\frac{\text{Max} V_{\text{in}}}{V_A}$ and the Max Rate Time on different half-widths ($r_0$) of the EMF. The right panel depicts the dependence of the Max Temperature and the Height on different half-widths ($r_0$) of the EMF. The length, temperature, time, and length units are $301.6 \text{ km}$, $10,000 \text{ K}$, $33.1 \text{ s}$, and $301.6 \text{ km}$, respectively.

Figure 10. This schematic shows the basic physical processes of the two types of simulated microflares, i.e., microflares with chromospheric or coronal origins. (A color version of this figure is available in the online journal.)

those with coronal chromospheric origin, can be illustrated by the schematic in Figure 10. According to our simulations, small microflares with weak Doppler velocities and only Hα emissions are most likely due to reconnection in the lower solar atmosphere (for instance, the chromosphere or photosphere). However, big microflares with obvious Doppler velocities and
emissions in both the low and high temperature wavelengths ($\text{Hz}$ and EUV/SXR) originate from reconnection in the corona. This model is similar to the unified model of Shibata (1996), Shibata et al. (2007), and Chen et al. (1999) in the sense that the different behaviors of the eruptions are determined by the different heights of the magnetic reconnection. According to the results in this paper, the strength and the size of the EMF determine whether the microflare is of a coronal or a chromospheric origin when the EMF emerges from the center of a supergranule, probably dragged by the convective upflow. Of course, EMFs may also appear at other sites of a supergranule, which would also affect the height of the reconnection site.

As mentioned in Section 3.1, we did find some plasmoids in our simulations. Figure 11 shows the current density distribution of a small region $[-35, 0] \times [10, 45]$ in Figure 3. From both the temperature and current density distributions (Figures 3 and 11), we can see that a plasmoid is formed at the center of the current sheet. The ejection of the plasmoid can temporarily increase the magnetic reconnection rate (Shibata et al. 1995; Shibata 1996; Shen et al. 2011). In the reconnection process, the plasmoids are generated one by one; therefore, we can see some oscillations of the reconnection rate in Figure 4. However, for the chromospheric case, the Alfvén speed is much smaller than in the coronal case, which leads to a smaller magnetic Reynolds number. Thus, similar to the results presented by Shen et al. (2011), there is no plasmoid produced in the chromospheric case. The corresponding reconnection rate in Figure 6 also shows less oscillations. In the coronal case, because the reconnection occurs in the corona, both the hot jets and the cold surges are ejected. There is no doubt that the EUV/SXR emissions from the hot jet will be observed first, whereas $\text{Hz}/\text{Ca}$ bright points and surges appear later. If we use the temperature response at a height of 1500 km, which is the $\text{Hz}$ formation height (Vernazza et al. 1981), the roughly calculated time delay between the EUV/SXR and $\text{Hz}$ emissions is about 3–5 minutes, which is comparable to the observations (Zhang et al. 2012). We must mention that the temperature enhancement and the time delay may be more realistic if we include the thermal conduction, radiation, the high-energy particles (nonthermal electrons and ions), and the effect of the partially ionized plasma. For simplicity as well as to see how the emerging fluxes with different sizes and strengths influence the reconnection process, we kept the threshold of the current density $j_c$ constant in the anomalous resistivity in Equation (8). In reality, the threshold of $j_c$ in the chromosphere might be higher because of its high plasma density. To see how $j_c$ would change the reconnection process, we performed other numerical experiments with higher $j_c$ in the chromosphere, and we found that the results are basically the same; the main difference is that the commencement of the reconnection process is only postponed for the case of the current sheet located at the chromosphere, as demonstrated by Yokoyama & Shibata (1994).

In summary, we successfully simulated the coronal and chromospheric microflares by using MHD simulations with a canopy-type magnetic configuration. The different observational behaviors between the coronal and chromospheric microflares are due to the height of the magnetic reconnection, which is determined by the size of the emerging flux, i.e., smaller emerging fluxes produce chromospheric microflares and larger emerging fluxes produce coronal microflares. Correspondingly, the resulting microflares have different sizes. In microflares with a coronal origin, the sizes of the simulated microflares are 11,000–16,000 km, i.e., $15^\circ$–$22^\circ$, which correspond to big microflares. We find a hot jet ($\sim 1.8 \times 10^6$ K for the typical case), which probably relates to the observational EUV/SXR jet, and a cold jet ($\sim 10^4$ K for the typical case), which corresponds to the observational $\text{H}\alpha$/Ca surge or brightening. Some of the plasmoids generated at the center of the current sheet are ejected, which can increase the magnetic reconnection rate. In microflares with a chromospheric origin, the sizes of the microflares are $4200$–$4500$ km, i.e., $\sim 6^\circ$, which correspond to relatively small microflares. Since the reconnection process occurs in the chromosphere, only the $\text{Hz}$/Ca brightenings show up in this case, and no significant SXR brightening or plasmoids are produced in this case. The parameter survey qualitatively shows that the size and strength of the EMF are the key parameters that determine the height of the reconnection X-point. For some typical values, we can achieve the reconnection process either in the corona or in the chromosphere. The parameter $\eta_{\text{max}}$ has little effect on the final results.

This paper deals with the microflares related to EMFs. It should be mentioned that some microflares are due to other mechanisms, e.g., the reconnection between the MMFs (Harvey & Harvey 1973) and the pre-existing magnetic field. Priest et al. (1994) proposed a canceling magnetic features model. Browning et al. (2008) performed three-dimensional MHD simulations to study the role of kink instability in the magnetic energy release process, which may also lead to microflares. It is also noted that the simulated emerging flux in this paper does not include the twisted field, which is required for the emerging flux to survive the subsurface convection (see Babcock 1961; Piddington 1975; Fan 2009, and references therein), e.g., Murray & Hood (2008) found that the tension force of the twisted tube plays a key role in determining whether the flux can successfully emerge to the upper solar atmosphere in three-dimensional numerical experiments. In the two-dimensional coronal simulations, the twisted field can be easily considered by including $B_z$, which would not greatly change the results. Moreover, a single-fluid model for the solar atmosphere was adopted in this paper for simplicity. In reality, the chromosphere, especially the lower part, is weakly ionized, and partial ionization should be considered, which would lead to strong anisotropic Cowling resistivity (Cowling...
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1957; Khodachenko et al. 2004; Leake & Arber 2006 and ambipolar diffusion (Vishniac & Lazarian 1999; Soler et al. 2009; Singh et al. 2011). Moreover, the ionization of the neutral atoms would consume a significant part of the released energy in magnetic reconnection (Chen et al. 2001; Jiang et al. 2010).

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