Similarity between the Molecular Loops in the Galactic Center and the Solar Chromospheric Arch Filaments

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(Received 2009 February 1; accepted 2009 May 27)

Abstract

We carried out two-dimensional magnetohydrodynamic simulations of the Galactic gas disk to show that the dense loop-like structures discovered by the Galactic center molecular cloud survey using the NANTEN 4-m telescope can be formed by a buoyant rise of magnetic loops due to the Parker instability. At the initial state, we assumed a gravitationally stratified disk consisting of a cool layer ($T \sim 10^3$ K), a warm layer ($T \sim 10^4$ K), and a hot layer ($T \sim 10^5$ K). The simulation box was a local part of the disk containing the equatorial plane. The gravitational field was approximated by that of a point mass at the Galactic center. The self-gravity, and the effects of the Galactic rotation were ignored. Numerical results indicate that the length of the magnetic loops emerging from the disk is determined by the scale height of the hot layer ($\sim 100$ pc at 1 kpc from the Galactic center). The loop length, velocity gradient along the loops, and large velocity dispersions at their foot points are consistent with the NANTEN observations. We also show that the loops become top-heavy when the curvature of the loop is sufficiently small, so that the rising loop accumulates the overlying gas faster than sliding it down along the loop. This mechanism is similar to that in the formation of solar chromospheric arch filaments. The molecular loops emerge from the low-temperature layer just like the dark filaments observed in the Hα image of the emerging flux region of the Sun.

Key words: Galaxy: magnetic loops, ISM — Sun: chromosphere — Sun: magnetic fields, magnetohydrodynamics

1. Introduction

Spiral galaxies have large-scale mean magnetic fields. It is widely accepted that the magnetic fields are amplified and maintained by the dynamo mechanism (e.g., Parker 1971). Mathewson and Ford (1970) showed by observations of optical polarization due to dust grains that the magnetic fields are nearly parallel to the galactic plane with wave-like patterns or loop-like structures, suggesting the emergence of a magnetic field via magnetic buoyancy (e.g., Parker 1966; Mouschovias et al. 1974; Blitz & Shu 1980; see also Tajima & Shibata 1997). A typical magnetic field strength in the Galactic plane is a few $\mu$G. The magnetic field strength increases with an increase in density toward the Galactic center. According to radio observations of the Galactic center and its vicinity, the magnetic field strength reaches a few mG in local regions near the Galactic center (Morris & Serabyn 1996).

The Galactic gas disk consists of a gas with three phases: namely, cold ($\sim 30$ K), warm ($\sim 8 \times 10^3$ K), and hot ($\sim 10^6$ K) components (McKee & Ostriker 1977). The cold component corresponds to molecular clouds observed with radio molecular lines (such as CO lines), while the warm component was observed with the H1 line, and the hot component was observed with soft X-rays.

The Galactic molecular cloud survey by the NANTEN 4-m telescope (Mizuno & Fukui 2004) found two dense gas features having a loop-like shape with a length of several hundred pc and width of $\sim 30$ pc within 1 kpc from the Galactic center (Fukui et al. 2006). Fukui et al. (2006) discovered large gradients in the line-of-sight velocity ($\sim 30$ km s$^{-1}$) along the molecular loops. Moreover, a large velocity dispersion ($\sim 30$–50 km s$^{-1}$) was observed near the foot-points of the...
loops. Since the loops have a total mass of $\sim 10^5 M_\odot$, the kinetic energy of a loop is estimated to be $\sim 10^{51}$ erg.

This energy is too large to be explained by a single supernova explosion. Moreover, the velocity distribution is distinct from that of an expanding shell. These features can be explained by the magnetic loops buoyantly rising due to the Parker instability (Parker 1966).

Matsumoto et al. (1988) carried out two-dimensional (2D) magnetohydrodynamic (MHD) simulations of the Parker instability. They found that dense regions are formed in the valley of magnetic loops in the nonlinear stage of the Parker instability, and that shock waves are formed at the footpoints of the rising magnetic loops where the gas infalling along the magnetic field lines collides with dense gas near the equatorial plane of the disk. Shibata and Matsumoto (1991) applied the results of the MHD simulations to the formation of molecular clouds in the Orion region. Subsequently, Kamaya et al. (1996) studied the triggering of the Parker instability by supernova explosions. Basu, Mouschovias, and Paleologou (1997) investigated the effect of the Parker instability on the structure of the interstellar medium, and Kim et al. (2000) applied it to galactic disks. Machida, Hayashi, and Matsumoto (2000) reported the results of three-dimensional (3D) global MHD simulations of the Parker instability in a differentially-rotating disk. They found that magnetic loops emerging from the disk form a structured corona above the disk.

Fukui et al. (2006) presented the results of 2D MHD simulations of local volume of the Galactic disk. They suggested that the downflows can be the origin of the violent motion and extensive heating of the molecular gas in the Galactic center. In their simulations, however, the cold component of the Galactic gas that corresponds to molecular clouds was not taken into account, and the mechanism of the formation of dense loop-like structures was not clear.

In the solar atmosphere, arch filaments were observed in the $H\alpha$ images of the chromosphere. Isobe et al. (2006) found that dense filaments similar to $H\alpha$ arch filaments are formed in the emerging flux. These filamentary structures are cool and dense above the chromosphere. This configuration in which the dense regions exist around the top of the emerging loops is a typical feature of the emerging loops, and is a good approximation to what will be observed as dark features in the $H\alpha$ image. They clarified the reason why the emerging loop becomes top-heavy on the basis of the results of 3D and 2D MHD simulations of the emergence of the magnetic loops from the convection zone, through the chromosphere, to the corona. The dense gas accumulated around the top of the rising loops fragments into filaments by the Rayleigh–Taylor instability, and slides down along the magnetic field lines. These filaments correspond to the arch filaments observed in the emerging flux regions of the Sun.

Fig. 1. Schematic picture of the simulation model and simulation box.
gravitational potential by a point mass, and gives a simple but exact model for an accretion disk rotating around the point mass. In the case of galaxies, this is not an exact model, but at least the region of the equator can be approximated by this model, because the gravitational acceleration near the disk in galaxies is approximately proportional to the height from the equatorial plane of the disk. The radial component of gravity is assumed to be equal to the centrifugal force due to the rotation of the disk. The effects of the rotation are discussed in subsection 4.4.

Equations (1)–(4) are rendered dimensionless by using normalizing constants $r_0$, $C_{S0}$, and $\rho_0$, where $C_{S0}$ is the sound speed in the mid-temperature region and $\rho_0$ is the unperturbed density at the equatorial plane. When the numerical results are compared with observations, we have used units of length, velocity, and time in the simulation to be $r_0 = 1$ kpc, $C_{S0} = 18$ km s$^{-1}$, and $t_0 = r_0/C_{S0} = 5.6 \times 10^7$ yr, respectively. The unit temperature is $T_0 = \mu C_{S0}^2/(\gamma R) = 2 \times 10^4$ K, where $\mu$ and $R$ are the mean molecular weight and gas constant, respectively. The normalized units are summarized in table 1.

We also introduce a non-dimensional parameter $\varepsilon = \sqrt{\gamma (V_K^2/C_{S0}^2)}$, where $V_K = (GM/r_0)^{1/2}$ is the Keplerian velocity at radius $r_0$. In this simulation, we have adapted $\varepsilon = 130$ and $V_K = 200$ km s$^{-1}$ for this parameter, because the sound speed in the low-temperature region is about $5.6$ km s$^{-1}$, whose value is close to the observed velocity dispersion ($5$–$9$ km s$^{-1}$) of the main gas components (Boulares & Cox 1990). However, this sound speed is larger than that of the molecular gas.

2.2. Initial State

The initial state is assumed to be in magnetohydrostatic equilibrium. The gas layer is initially composed of three layers: a cool equatorial layer ($T = T_c$, $|z| < z_1$), a warm (mid-temperature) layer ($T = T_m$, $z_1 \leq |z| \leq z_2$), and a hot Galactic halo ($T = T_h$, $|z| > z_2$). The initial distribution of temperature is assumed to be

$$T(z) = T_c + (T_m - T_c) \left\{ \frac{1}{2} \left( \tanh \left( \frac{|z| - z_1}{w_1} \right) + 1 \right) \right\}$$

$$+ (T_h - T_m) \left\{ \frac{1}{2} \left( \tanh \left( \frac{|z| - z_2}{w_1} \right) + 1 \right) \right\}.$$  (6)

In this study, we took $T_c = 0.1 T_0$, $T_m = T_0$, $T_h = 10 T_0$, $w_1 = 0.015 r_0$, $z_1 = 0.12 r_0$, and $z_2 = 0.24 r_0$. Although these values are not realistic for the Galactic gas disk, they would be acceptable for our first attempt to study the effects of the cold component of the Galactic gas and the mechanism for the formation of dense loop-like structures. Numerical results do not depend much on the temperature of the hot Galactic halo (Kamaya et al. 1997). Further discussion on the dependence of the numerical results on the disk temperature are given in subsection 4.5.

We assume that the magnetic field is initially parallel to the equatorial plane; $B = [B_z(z), 0, 0]$, and is localized in the cool equatorial region ($|z| < z_1$). The magnetic field strength is determined by introducing the plasma beta (the ratio of the gas pressure to the magnetic pressure, hereafter denoted by $\beta$) as

$$B_z(z) = \sqrt{\frac{8\pi P(z)}{\beta(z)}},$$  (7)

where

$$\frac{1}{\beta(z)} = 1 + \frac{1}{\beta_0} \left\{ 1 - \frac{1}{2} \left( \tanh \left( \frac{|z| - z_t}{w_1} \right) + 1 \right) \right\}.$$  (8)

Here, $\beta_0$ is the plasma $\beta$ at the Galactic plane, and $z_t$ is the half thickness of the magnetic flux sheet. The initial density and gas-pressure distributions are calculated numerically by solving the equation of magnetohydrostatic equilibrium,

$$\frac{d}{dz} \left[ P(z) + \frac{B_z^2(z)}{8\pi} \right] + \rho(z) g(z) = 0.$$  (9)

We have studied the evolution for five models, whose parameters are summarized in table 2. The distributions of the initial temperature, $T$, density, $\rho$, gas pressure, $P$, and magnetic pressure, $B_z^2/(8\pi)$ are shown in figure 2a. Figure 2b shows the profile of the initial local pressure scale height, $\Lambda(z) = C_S(z)^2/[\gamma g(z)]$, and the pressure scale height including the magnetic field, $\Lambda_{B}(z) = [1 + \beta(z)^{-1}] \Lambda(z)$, for model B ($\beta_0 = 1$), which we call the fiducial model hereafter.

2.3. Stability of the Equilibrium Model

The equilibrium state that we described in subsection 2.2 is unstable against the Parker instability, which is in an isothermal gas and has a linear maximum growth rate at a finite
Fig. 2. (a) Distribution of the initial density (dashed curve), gas pressure (dotted curve), magnetic pressure (solid curve), and temperature (dash-dotted curve), for the fiducial model (model B). (b) Distribution of the initial local pressure scale height $\Lambda$ (solid curve) and modified scale height $\Lambda_B = [1 + \beta(z)^{-1}]\Lambda$ (dashed curve) for the fiducial model (model B).

Fig. 3. (a) Linear growth rate of the Parker instability for the unperturbed states with $\beta_0 = 0.5, 1, 2, 4,$ and $10$ as a function of the wave number, $k_x$. (b) Dependence of the maximum growth rates on the initial plasma beta. The solid line shows a line where $i\omega \propto \beta_0^{-1/2}$. The units of the growth rate are $C_{S0}/r_0 = 1/r_0$.

Figure 3a shows the linear growth rate ($i\omega$) of the fundamental mode (Horiuchi et al. 1988) of the Parker instability as a function of the horizontal wave number, $k_x$, for five cases: $\beta_0 = 0.5, 1, 2, 4,$ and $10$ when $z_f/r_0 = 0.08$ and $k_y = 0$. Figure 3b shows that the maximum growth rate is inversely proportional to the square root of $\beta$ ($i\omega \propto \beta_0^{-1/2}$). This result is consistent with that of a linear analysis of the Parker instability in uniform gravitational fields (Parker 1966). The most unstable wavelength ($\lambda_{max}$) is $\lambda_{max} \approx (\pi/8)r_0$ for model B ($\beta_0 = 1$). It is noted from figures 2b and 3 that $\lambda_{max}$ is nearly 10-times the local pressure scale height of the mid-temperature wavelength, $\lambda_{Parker} \sim 10\Lambda-20\Lambda$, where $\Lambda$ is the scale height (Parker 1966). This is because small-wavelength modes are stabilized by a magnetic tension force. We analyzed the linear stability of the initial model with a normal-mode method similar to that of Horiuchi et al. (1988). The linearized equations are the same as those in Horiuchi et al. (1988) and Kamaya et al. (1997). We consider the growth of a small perturbation that has a functional form of $\delta W = \exp(i\omega t + ik_xx + ik_yy)$, where $W$ is a physical quantity ($p, P, v_z, B_z$) and $\delta W$ is its perturbation. The eigenvalues ($\omega$) and the eigenfunctions were calculated numerically.
Fig. 4. Numerical results for the fiducial model (model B); density in a logarithmic scale (colors), magnetic field lines (solid curves), and velocity field (vectors) for $t/t_0 = 0.0, 0.5, 1.0, 1.2, 1.4, \text{ and } 1.6$.

region in $0.12 \leq z/r_0 \leq 0.24$ ($\lambda_{\text{max}} \sim 10 \Lambda \sim 400 \text{ pc}$), and is much larger than that of the low-temperature cool region around $z/r_0 \sim 0.1$.

2.4. Boundary Conditions and Numerical Method

We assumed free boundaries at $z = Z_{\text{min}}$ and $z = Z_{\text{max}}$ such that waves transmit freely by setting the $z$-derivatives of all the variables to vanish, and imposed periodic boundary conditions at $x = X_{\text{min}}$ and $x = X_{\text{max}}$.

In order to trigger the Parker instability, small-velocity perturbations of the form

$$v_z = A C_{S0} \cos \left( \frac{2 \pi x}{\lambda_c} \right)$$

are given initially within a finite horizontal domain ($|x| \leq \lambda_c/2$) on the magnetic flux sheet, where the perturbation wavelength, $\lambda_c$, is close to that of the most unstable wavelength for the Parker instability. Here, $A (= 0.01)$ is the maximum value of $v_z/C_{S0}$ in the initial perturbation.

The numerical scheme that we used is the Rational CIP (Cubic interpolated profile) method (Yabe & Aoki 1991; Xiao et al. 1996) combined with the MOC-CT method (Evans & Hawley 1988; Stone & Norman 1992). The magnetic induction
equation was solved by the MOC-CT, and the other equations were solved by the CIP (Kudoh et al. 1998, 1999).

The size of the simulation box was $(X_{\text{max}} - X_{\text{min}}, Z_{\text{max}} - Z_{\text{min}}) = (4r_0, 4r_0) = (19H, 19H)$, where $H$ is the scale height at the point where the gravity is maximum $[H = C_2^2(\gamma g_{\text{max}}) = 0.21r_0]$. The grid sizes were $\Delta x = 2.5 \times 10^{-3}r_0$ for $|x| \leq 0.8r_0$, $\Delta z = 2.5 \times 10^{-3}r_0$ for $|z| \leq 0.8r_0$, and they slowly increased for $|x| > 0.8r_0$, or $|z| > 0.8r_0$ by an increment of 5% at each grid (e.g., $|\Delta x_{j+1}| = 1.05|\Delta x_j|$). The number of grid points was $(N_x, N_z) = (472, 472)$.

### 3. Numerical Results

Figure 4 shows the time evolution of the fiducial model. The solid curves depict magnetic field lines. Color shows the density distribution. Arrows show velocity vectors. The overall evolution agrees with that of Matsumoto et al. (1988). That is to say, as the instability grows, the magnetic field lines bend across the equatorial plane. As the gas slides down along the undulating magnetic field lines, the rarefied regions buoyantly rise, and form magnetic loops in the later (nonlinear) phase. In the valleys of the magnetic loops, dense spur-like structures are created almost perpendicular to the Galactic plane. At the top of the emerging loops, dense shell-like structures are formed (figure 5a). These structures were not recognized in previous simulations (e.g., Matsumoto et al. 1988), partly because they assumed an isothermal atmosphere without a steep density gradient at the disk–halo interface, and partly because of the lower numerical resolution. Since the downflow speed exceeds the local sound speed, strong shock waves are formed at the magnetic loop footpoints. Figure 5b shows that the shock waves heat the cool gas at around $z/r_0 \sim 0.1$.

At the final phase ($t/t_0 = 1.6$), the expansion of the magnetic loops is stalled because the driving forces diminish at the top of the loop. Figure 6 shows the vertical distribution of the magnetic pressure, gas pressure, $\beta$, and density at the midpoint of the emerging loop ($x/r_0 = 0.25$). From this figure, it is clear that the magnetic pressure and gas pressure gradients nearly balance at the top of the emerging loop. This magnetohydrostatic state is attained because the Parker instability is stabilized in the hot layer where the local pressure scale height becomes larger than the length of the magnetic loop.

Figure 7 shows snapshots of the density distribution for all models ($\beta_0 = 0.5, 1, 2, 4,$ and $10$) at the stage when the top of the magnetic loops enters the hot region ($z/r_0 > 0.24$). The time scale for the loop emergence is shorter for lower $\beta$, consistent with the results of a linear analysis presented in subsection 2.3. Figure 8 shows the vertical distributions of the density at the midpoint of the emerging loop. The density increases with height at the top of the magnetic loops in all models. In the model with $\beta_0 = 10$, the emerging loop stalls at a lower height ($z/r_0 < 0.3$) because the released magnetic energy is small.
Fig. 7. Numerical results for all models ($\beta_0 = 0.5, 1, 2, 4, \text{and} 10$); density on a logarithmic scale (colors), magnetic field lines (solid curves), and velocity field vectors.
4. Discussion

4.1. Formation of Dense Loop Structure

Let us discuss why the emerging loop becomes top-heavy. Figure 9 shows the magnetic field lines at $t/t_0 = 1.4$ (solid curves) and $t/t_0 = 0.9$ (broken curves) for the fiducial model (model B). These field lines are iso-contours of the $y$-component of the vector potential. The magnetic field lines comoving with the plasma in ideal MHD can be identified by the value of the vector potential. Since the frozen-in condition is assumed and the numerical diffusion is negligibly small, the field lines at $t/t_0 = 0.9$ (broken curves) have moved with gas to the corresponding field lines at $t/t_0 = 1.4$ (solid curves). As the magnetic loops rise, they can accumulate the gas above the loop when the loop top is flat. Figure 9 shows that the loop shape at $t/t_0 = 0.9$ is favorable for mass accumulation around the loop top. At $t/t_0 = 1.4$, since the curvature increases, the mass accumulated around the loop top slides down along the magnetic field lines. Flat top loops are formed when the loop length is much longer than the local pressure scale height. In our model atmosphere, since the local pressure scale height decreases with height in $0 < z/r_0 < 0.1$ and increases with height in $0.1 < z/r_0 < 0.25$ (see figure 2b), the loop top tends to be flat when its height is $z/r_0 \sim 0.1$. A dense shell around the loop top is formed when $0.1 < z/r_0 < 0.2$, and the mass drains as the loop rises.

In order to quantitatively evaluate the density enhancement at the top of the emerging loop, we adopt the same method of analysis as Isobe et al. (2006). The velocity vector $\mathbf{v}$ can be divided into two components:

$$\mathbf{v} = v_\perp + v_\parallel,$$

where $v_\perp$ and

$$v_\parallel = \frac{\mathbf{v} \cdot \mathbf{B}}{|B|^2} \mathbf{B}$$

are the velocity components perpendicular and parallel to the magnetic field, respectively.

Figure 10 shows the distribution of $\mathbf{v}$ and $\nabla \cdot \mathbf{v}$ (figures 10a and 10b), $v_\perp$ and $\nabla \cdot v_\perp$ (figures 10c and 10d), and $v_\parallel$ and $\nabla \cdot v_\parallel$ (figures 10e and 10f) at $t/t_0 = 1.2$ and $t/t_0 = 1.3$ for the fiducial model. The distributions of $\nabla \cdot v_\perp$ and $\nabla \cdot v_\parallel$ show that the density inside the emerging magnetic loops keeps decreasing. However, $\nabla \cdot v_\parallel$ at the top of the emerging loop are negative. Therefore, at least for the present parameters, it can be concluded that the density increases in the top region of the emerging loop. Figure 11 shows the distribution of the density near the loop top at $t/t_0 = 1.1$ and $t/t_0 = 1.2$ for the fiducial model. The clump of gas near the loop top survived due to a small curvature in the low-temperature layer, and is later compressed in the high-temperature layer.

4.2. Comparison with Observations

Fukui et al. (2006) found two dense gas features having a loop-like shape with a length of several hundred pc within ~1 kpc from the Galactic center. Figure 3a shows that the most unstable wavelength is likely to be 350–500 pc. Our numerical simulation produced magnetic loops whose wavelength is ~400 pc and their height is ~350 pc. In the observed molecular loops, the line-of-sight velocity along the loop changes linearly with the arc-length, and has large gradients (~30 km s$^{-1}$). Figure 12 shows the velocity components...
along the outermost magnetic field line at $t/t_0 = 1.4$ for the fiducial model. The gradient of the downflow velocity given in figure 12c is $\sim 20-30 \text{ km s}^{-1}$. Since the downflow speed exceeds the local sound speed, strong shock waves are formed at the magnetic loop footpoints. Such shocks can be the origin of the observed large velocity dispersions near the footpoints of the Galactic center molecular loops.

### 4.3. Estimation of the Kinetic Energy of the Emerging Loop

Fukui et al. (2006) estimated the kinetic energy of the molecular loop to be $\sim 10^{51} \text{ erg}$. In our simulation, the kinetic energy flux, $F_k$, carried by the downflow is

$$F_k \approx \frac{1}{2} \rho_{\text{df}} v_{\text{df}}^3 \sim 10^{-5} \text{ erg cm}^{-2} \text{ s}^{-1},$$

(13)
where $\rho_{df}$ is assumed to be $10^{-24}$ g cm$^{-3}$ and $v_{df}$ is found to be $\sim 30$ km s$^{-1}$ according to a numerical result for the fiducial model. The numerical results indicate that the loop width in the $z$-direction is about 30 pc ($\sim 10^{20}$ cm) and the downflow continues for about $10^{15}$ s $\sim 3 \times 10^7$ yr. When the loop thickness in the $y$-direction is 100 pc, the total kinetic energy of the downflow toward the both footpoints is $\sim 10^{51}$ erg. This energy is comparable to that estimated from the observation.

4.4. Effects of Cooling and Rotation

In this study, we assumed that the gas layer at the initial state is composed of the cool equatorial, warm, and hot layers. We assumed adiabatic gas, but in the interstellar gas, cooling and heating play essential roles in the formation of molecular clouds (Field 1965). Kosiński and Hanasz (2006, 2007) investigated the Parker instability coupled with thermal processes (cooling and heating). They found that the Parker instability can trigger a thermal instability that form dense clouds in the valleys of the magnetic loops. In subsequent papers, we would like to include gas cooling and heating effects.

We should consider the effect of the rotation of the disk to construct a more realistic model. In rotating disks, the Parker instability is slightly suppressed by Coriolis forces (Chou et al. 1997). Hanasz, Otmianowska-Mazur, and Lesch (2002) presented the results of resistive 3D MHD simulations of a local part of the disk including the contribution of the Coriolis force. They demonstrated that the Parker instability, the twisting of the loops, and magnetic reconnection lead to the formation of helically twisted magnetic flux tubes. It is worth noting other 3D effects, such as the growth of an interchange mode (e.g., Nozawa 2005). Isobe et al. (2006) showed that dense shells at the top of the magnetic loops fragment into filaments due to the growth of the Rayleigh–Taylor instability. In order to include this effect, we need to carry out 3D MHD simulations.

In Galactic gas disks, the magneto-rotational instability (MRI: Balbus & Hawley 1991) grows and drives magnetic turbulence inside the disk. In the nonlinear stage, MRI drives magnetic turbulence inside the disk. The effects of stochastic magnetic fields on the growth of the Parker instability was reported by Parker and Jokipii (2000). The results of global 3D MHD simulations of the Galactic center gas disk are reported by Machida et al. (2009).

4.5. Dependence on the Initial Disk Temperature

We studied the dependence of the numerical results on the initial disk temperature. Figure 13 shows the density distribution at the stage when the top of the magnetic loops reaches $Z/r_0 \sim 0.35$ for models with a lower initial disk temperature ($T_c = 0.05T_0$) and a higher initial disk temperature ($T_c = 0.2T_0$). The other parameters are the same as the fiducial model (model B). The numerical results indicate that the equatorial dense region at this stage is thin for the cool disk ($T_c = 0.05T_0$) and thick for the warm disk ($T_c = 0.2T_0$). Although the length of the loops near the equator is smaller for the cooler disk, the length of the loops emerging in the halo is almost the same ($\sim 400$ pc). This is because the most unstable wavelength of the Parker instability is determined by the local pressure scale height.

4.6. Other Effects

Finally, we briefly mention about other effects, such as cosmic rays, self-gravity, and the presence of a spiral arm of the Galactic gas disk. In galactic disks, since the cosmic-ray

Fig. 11. Distribution of the logarithmic density (solid curve) and pressure (broken curve) near the loop top at $x/r_0 = 0.25$ for the fiducial model (model B) at $t/t_0 = 1.1$ (left) and $t/t_0 = 1.2$ (right).
pressure is comparable to the gas pressure, cosmic rays enhance the growth rate of the Parker instability (e.g., Parker 1966; Hanasz & Lesch 2000, 2003; Kuwabara et al. 2004). The effects of the cosmic rays may be more important in the Galactic center where strong activities can produce high-energy particles.

In this paper, we neglected the self-gravity of the gas. When the surface density of the gas disk is large enough, the Parker–Jeans instability grows (e.g., Elmegreen 1982; Nakamura et al. 1991; Kim et al. 2002; Lee et al. 2004). The Jeans instability has a larger growth rate than does the Parker instability when either the magnetic field is weak or the wavelength of the perturbation is long.

Franco et al. (2002) showed by 3D MHD simulations that the Parker instability creates massive clouds inside the spiral arm of the Galactic gas disk, and that dense gas accumulated around the equatorial plane forms a corrugated structure. The distribution of H I gas and dust below the molecular loops found in the Galactic center (Torii et al. 2009) is consistent with the simulation by Franco et al. (2002). The density distribution in the nonlinear stage of our simulation also shows such corrugated structures.
5. Summary

In the present work, we carried out 2D MHD simulations of the Galactic center gas disk consisting of a low-temperature layer and a mid-temperature layer with an overlying hot halo. We found that the numerical results reproduce basic features in the Galactic center molecular loops, as observed with NANTEN, such as the loop length, the velocity gradient along the loops, and large velocity dispersions at the footpoints of the loops.

We also discussed the reason why the top of the emerging loop becomes over-dense. This is because the effective gravity along the magnetic field lines decreases when the curvature radius of the magnetic field lines increases in the low-temperature layer. The gas that survives near the loop top due to its small curvature is compressed when the loop top enters the high-temperature layer.

Finally, we suggest that the Galactic center molecular loops are analogous to the arch filaments in the solar chromosphere. The molecular loops emerge from the low-temperature layer just like the dark filaments observed in H$\alpha$ images of the emerging flux region of the Sun.

We are grateful to T. Sakurai, T. Kudoh, and D. Shiota for useful comments and discussion. Numerical computations were carried out on the general-purpose PC farm at Center for Computational Astrophysics, CfCA, of National Astronomical Observatory of Japan (P.I. KT). This work is financially supported in part by a Grant-in-Aid for Scientific Research (KAKENHI) from Japan Society for the Promotion of Science (JSPS) (No. 20244014). This work is also carried out by the joint research program of the Solar-Terrestrial Environment Laboratory, Nagoya University.

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