THE MASS OF THE FIRST STARS

Hajime Susa
Department of Physics, Konan University, Okamoto, Kobe, Japan; susa@konan-u.ac.jp
Received 2013 January 27; accepted 2013 July 5; published 2013 August 6

ABSTRACT

We perform a three-dimensional radiation hydrodynamics simulation to investigate the formation of the first stars from the initial collapse of a primordial gas cloud to the formation and growth of protostars. The simulation is integrated until \( \sim 0.1 \) Myr after the formation of the primary protostar, by which time the protostars have already settled onto the main sequence. This work represents the first attempt at simulating the first episodes of star formation, taking into account the ultraviolet radiative feedback effect from multiple protostars as well as the three-dimensional effects of the fragmentation of the accretion disk. We find that the mass accretion onto Population III protostars is significantly suppressed by their radiative feedback. As a result, we find five stars formed in this particular simulation. The final masses of the stars are \( \lesssim 60 \, M_\odot \), including a star of 4.4 \( M_\odot \). Formation of such a star hints at the existence of even lower-mass stars that would live today.

Key words: dark ages, reionization, first stars – hydrodynamics – radiative transfer – stars: formation – stars: Population III

Online-only material: color figures

1. INTRODUCTION

The formation of the first stars has been investigated intensively in the last decade, mainly from theoretical aspects. Following the theoretical predictions, the first stars form in minihalos of masses \( \sim 10^5 - 10^6 \, M_\odot \) (Haiman et al. 1996; Tegmark et al. 1997; Nishi & Susa 1999; Fuller & Couchman 2000; Abel et al. 2002; Bromm et al. 2002; Yoshida et al. 2003).

An ingredient of the formation of the first stars is primordial gas, which does not contain heavy elements or cosmic dust. Because of the lack of these efficient coolants, the primordial gas cools inefficiently, especially at low temperatures (\( T \lesssim 10^4 \) K). Therefore, the gas is kept relatively warm (\( \sim 10^3 \) K) while it collapses to form stars, in contrast with interstellar gas, whose temperature is \( \sim 10 \) K during present-day star formation for \( n_H \lesssim 10^{10} \, \text{cm}^{-3} \) (e.g., Omukai 2000). As a result, the gravitationally collapsing primordial clouds are very massive (\( \sim 1000 \, M_\odot \)), since they have to be more massive than the Jeans mass, which is proportional to \( T^{3/2} \). In addition, the formation of such a massive prestellar core leads to a huge mass accretion rate onto the protostar in the later mass accretion phase (e.g., Omukai & Nishi 1998; Bromm et al. 1999; Abel et al. 2000, 2002; Nakamura & Umemura 2001; Yoshida et al. 2006). Following the theoretical evidence, most of the first stars are expected to have once been very massive (\( \gtrsim 100 \, M_\odot \)).

On the other hand, studies of the mass accretion phase have advanced recently, revealing that the first star formation processes seem to be more complicated than expected before (Turk et al. 2009; Clark et al. 2011a, 2011b; Smith et al. 2011; Greif et al. 2011, 2012). These studies found that a heavy disk formed around the primary protostar because of the angular momentum of the prestellar core gained by tidal interactions with other cosmological overdense regions. The heavy disk fragments into small pieces since it is gravitationally unstable. As a result, a star cluster can be formed instead of a single, very massive star.

These results seem to be robust until the primary protostar grows to \( \gtrsim 20 \, M_\odot \). After the mass of the protostar exceeds \( \sim 20 \, M_\odot \), significant ultraviolet radiation flux will be emitted from the protostar (Omukai & Palla 2003; Hosokawa & Omukai 2009; Hosokawa et al. 2011). The ultraviolet radiation from the protostar significantly affects the later evolution of the system.

Hosokawa et al. (2011) have addressed this feedback effect directly by two-dimensional (2D) radiation hydrodynamics simulations. They found that the accretion disk around the primary protostar is photoevaporated due to radiative feedback, followed by the rapid decline of the mass accretion rate onto the protostar. The final mass of the protostar in their simulation is 43 \( M_\odot \).

Stacy et al. (2012) also tried to assess the final mass of the first stars in their three-dimensional cosmological calculations, where their methodology included the fragmentation of the disk as well. They found that the ultraviolet radiative feedback strongly suppresses mass accretion onto the protostars. However, the integrated physical time is \( \sim 5000 \) yr, which is too short to predict the final mass of the first stars since it will take \( \sim 0.1 \) Myr for the protostar to settle onto the main sequence (e.g., Hosokawa & Omukai 2009).

In this paper, we report the results of three-dimensional radiation hydrodynamics simulations on the formation of first stars that follow the evolution of the system for 0.1 Myr after the formation of the primary protostar. We take into consideration three-dimensional effects, as well as radiative feedback from the protostars.

2. NUMERICAL SIMULATIONS

We study the formation of the first stars using radiation hydrodynamics simulations. We employ a Bonner–Ebert sphere with \( n_H = 10^4 \, \text{cm}^{-3} \), \( T = 200 \) K as initial conditions of the simulation. The initial conditions are motivated by cosmological simulations (Abel et al. 2002; Yoshida et al. 2003) in which such clouds are found in minihalos of masses \( 10^5 - 10^6 \, M_\odot \) in the “loitering” phase. The loitering phase corresponds to the epoch when the cloud becomes quasi-static, because \( H_2 \) line cooling at \( n_H \gtrsim 10^4 \, \text{cm}^{-3} \) becomes less efficient than that at \( n_H \lesssim 10^4 \, \text{cm}^{-3} \). In order to make the gas cloud slightly gravitationally unstable, we increase the density by 20%. As a result, the total mass of the cloud is 2600 \( M_\odot \). We also impose uniform rotation on the gas sphere with \( \Omega = 2 \times 10^{-14} \, \text{rad s}^{-1} \).
The angular velocity of this value results in a very similar specific angular momentum distribution to that found in the cosmological simulations of Yoshida et al. (2006) at the final stage of the runaway collapse phase.

We use a radiation smoothed particle hydrodynamics (RSPH; Susa & Umemura 2004; Susa 2006) code in order to solve the equations of hydrodynamics, the non-equilibrium primordial chemistry of six species ($e^-$, $H^+$, $H$, $H_2$, $H^{-}$, and $H_3^+$), and the radiative transfer of ultraviolet photons. The reactions included in the code are listed in Table 1. Transfer of ultraviolet radiation is assessed by a ray-tracing scheme, in which neighbor smoothed particle hydrodynamics (SPH) particles are connected to make up the rays (Susa 2006). Then, we calculate the optical depth at the Lyman limit as well as the $H_2$ column density by the ray-tracing scheme. Using the optical depth at the Lyman limit, we can assess the correct photoionization/photoheating rate by integrating the spectrum before we start our simulation. We also remark that we employ an on-the-spot approximation for the ionizing photons. The $H_2$ column density is used to calculate the self-shielding function of Lyman–Werner (L–W) photons. We use an updated self-shielding function for the L–W radiative transfer (Wolcott-Green et al. 2011). $H_2$ photodissociation is also taken into account, based on the cross-section in Stancil (1994). $H^-$ radiative detachment is assessed using the fitting formula for the cross-section in Teegmark et al. (1997). We also assume that the radiation below the L–W band is optically thin.

We include the cooling processes of primordial gas such as $H/\mathrm{H}_2$ line cooling, $H_2$ formation heating/dissociation cooling, $H$ ionization/recombination cooling, Bremsstrahlung, optically thin $H^-$ cooling, and collision-induced emission (CIE) cooling.\(^1\)

---

We take into account the shielding of $H_2$ line emission, utilizing the shielding function proposed by Ripamonti & Abel (2004). In the present version of the RSPH code, we update the hydrodynamics gravity and the radiative transfer at the Courant time, while the energy equation and chemical reaction equations are integrated at smaller time steps.

The mass of an SPH particle in the present work is set to be $m_{\text{SPH}} = 4.96 \times 10^{-3} \, M_\odot$ and the number of neighbor particles is $N_{\text{neib}} = 50$, corresponding to a mass resolution of $M_{\text{res}} = 2 N_{\text{neib}} m_{\text{SPH}} = 0.496 \, M_\odot$ (Bate & Burkert 1997). This mass resolution is equivalent to the Jeans mass of $n_H = 10^{12} \, \text{cm}^{-3}$, $T = 300 \, \text{K}$, and comparable to that in Stacy et al. (2012).

In order to trace the evolution in the mass accretion phase, we employ sink particles. If the density at an SPH particle exceeds $n_{\text{sink}} = 5 \times 10^{13} \, \text{cm}^{-3}$, we change the SPH particle into a sink particle. In addition, if SPH particles fall within a sphere of radius $r_{\text{acc}} = 30 \, \text{AU}$ centered on a sink particle, and they are gravitationally bound to each other, these SPH particles are merged to the sink particle, conserving linear momentum and mass. The accretion radius $r_{\text{acc}}$ is again comparable to that employed in Stacy et al. (2012). We remark that sink–sink merging is not allowed in the present numerical experiment, since the radius of the protostar is less than ~1 AU (Hosokawa & Omukai 2009), which is much smaller than the employed accretion radius $r_{\text{acc}}$. We regard the mass of the sink particles as the mass of protostars. We also assume that the sink particles do not push surrounding SPH particles, i.e., the pressure forces from sink particles on the surrounding SPH particles are omitted. The recipe for the sink particles employed in the present work is known to overestimate the mass accretion rate (e.g., Bate et al. 1995; Bromm et al. 2002; Martel et al. 2006). In addition, the employed accretion radius $r_{\text{acc}}$ is 30 AU, which is much larger than the radius of protostars (Hosokawa & Omukai 2009). Thus, we have to keep in mind that the resultant mass of the formed sink particles would be larger than the actual mass of the first stars. We also remark that we cut the central spherical region with a radius of 0.6 pc out of the cloud, just after the formation of the first sink, in order to save computational time. The outer envelope of $r > 0.6 \, \text{pc}$ hardly affects the inner region before $10^5 \, \text{yr}$.

We turn on the sinks (stars) when they are created, although their initial masses are so small that very few ultraviolet photons are emitted initially. The mass accretion rates onto the sinks in the present simulation are obtained by averaging over $10^3 \, \text{yr}$ in order to avoid artificial discontinuities due to SPH discreteness. We feed this mass accretion rate into the protostellar evolution model to assess the stellar radius/effective temperature.\(^2\) The evolving luminosity and effective temperature of a protostar is obtained based on the calculations by Hosokawa & Omukai (2009). They calculated the evolution of protostars with given (fixed) mass accretion rates. On the other hand, we obtain the protostellar masses ($M$) and the mass accretion rates ($\dot{M}$) self-consistently from the hydrodynamics simulation. Then, we can assess the luminosities and effective temperatures by interpolating the data from Hosokawa & Omukai (2009) at every time step. These luminosities and temperatures are used to calculate the luminosities and blackbody spectra of the

---

\(^1\) CIE is added just for completeness, since it is important only at $n_H \gtrsim 10^{14} \, \text{cm}^{-3}$ (Yoshida et al. 2008).

\(^2\) We have to keep in mind the limitation of the present treatment of the protostellar model in which steady accretion is assumed (e.g., Stahler et al. 1986). More violent/clumpy mass accretion could maintain a cooler temperature of the protostar (Smith et al. 2012) for larger masses.

---

### Table 1

| Number | Chemical Reactions | Reference |
|--------|---------------------|-----------|
| 1      | $H^+ + e^- \rightarrow H + \gamma$ | SP        |
| 2      | $H + e^- \rightarrow H^+ + \gamma$ | GP        |
| 3      | $H + H^- \rightarrow H_2 + e^-$ | GP        |
| 4      | $3H \rightarrow H_2 + H^-$ | PSS       |
| 5      | $H + H_2 \rightarrow 3H$ | SK        |
| 6      | $2H + H_2 \rightarrow 2H + H_2$ | PSS       |
| 7      | $2H_2 \rightarrow 2H + H_2$ | SK        |
| 8      | $H + e^- \rightarrow H^+ + 2e^-$ | SK        |
| 9      | $2H + H^+ + e^- \rightarrow H_2 + H^+$ | PSS       |
| 10     | $H + H^+ \rightarrow H_2 + \gamma$ | GP        |
| 11     | $H_2^+ + H_2 \rightarrow H_2 + H^+$ | GP        |
| 12     | $H_2 + H^+ \rightarrow H_2^+ + H$ | GP        |
| 13     | $H^+ + H^- \rightarrow H_2^+ + e^-$ | GP        |
| 14     | $H_2^+ + e^- \rightarrow 2H$ | SK        |
| 15     | $H^+ + e^- \rightarrow H + 2e^-$ | SK        |
| 16     | $H^+ + H^- \rightarrow 2H$ | GP        |
| 17     | $H_2 + e^- \rightarrow H + H_2^+$ | GP        |
| 18     | $H_2 + e^- \rightarrow 2H_2^+$ | GP        |
| 19     | $H^+ + e^- \rightarrow H^- + H^+$ | SU        |
| 20     | $H_2^+ + e^- \rightarrow 2H$ | KBS+WHB   |
| 21     | $H_2^+ + e^- \rightarrow H + H^+$ | TG        |
| 22     | $H^- + \gamma \rightarrow H + e^-$ | ST        |

References. SP: Spitzer 1978; GP: Galli & Palla 1998; SK: Shapiro & Kang 1987; PSS: Palla et al. 1983; SU: Susa 2006; KBS: Keper et al. 1997; ST: Stancil 1994; WHB: Wolcott-Green et al. 2011; TG: Teegmark et al. 1997.
protostars. Hence, the protostellar evolution model is self-consistently fed to the radiation hydrodynamics calculations.

3. RESULTS

We perform radiative hydrodynamics simulations of the first star formation with radiative feedback. We also perform a run with no feedback for comparison.

3.1. Fragmentation of the Disk Around the Primary Protostar

We start the simulation from a rigidly rotating Bonner–Ebert sphere around the loitering phase. The cloud starts to collapse in a runaway fashion, i.e., the central density keeps growing while the outer part of the cloud is left in the envelope. As a result, a density profile of $\propto r^{-2.2}$ is built up during the runaway phase (Yoshida et al. 2006).

Eventually, the central density exceeds $n_{\text{sink}}$, and a sink particle is formed at the center of the cloud. The surrounding gas subsequently starts to accrete onto the sink particle. Since the gas has a significant amount of specific angular momentum, the accreting gas forms an accretion disk around the sink particle. The amount of specific angular momentum in the runaway phase is close to that of the similarity solution, which is approximately 0.5 times the value of Keplerian rotation at the Jeans radius, i.e., the core radius (Yoshida et al. 2006). Thus, the radius of the disk is 0.25 times smaller than the original radius in the runaway phase, since the centrifugal force is proportional to the square of the specific angular momentum.

After the formation of the small gas disk around the first sink, gas keeps accreting onto the disk. As a result, the mass and the radius of the disk increase. At the same time, the temperature of the disk decreases by radiative cooling. The left column of Figure 1 shows a face-on view of the disk column density at three epochs corresponding to 320 yr, 620 yr, and 860 yr after the formation of the first sink. The red crosses denote the positions of sink particles.

In the early epoch of the accretion phase, a smooth disk forms around the sink particle (top panel), followed by the formation
The Astrophysical Journal, 773:185 (8pp), 2013 August 20

Susa

Figure 2. Edge-on views of the gas distribution inside $r < 10^4$ AU (0.05 pc) at four time snapshots. Top row: from left to right, $t = -10$ yr, 1160 yr. Bottom row: $t = 5120$ yr, 100,250 yr. $t$ represents the time after the first sink formation. The color shows the H$_2$ fraction and the small white spheres represent the positions of the sink particles. The white arrow and the dashed curve in the bottom left panel denote the position of the shock front. (A color version of this figure is available in the online journal.)

of spiral arms (middle panel), and the fragmentation of the arms (bottom panel). We can also find a few high column-density peaks in the bottom panel. In fact, the next sink particles are born from these peaks within a few hundred years.

The right column of Figure 1 shows color contours of the Toomre $Q$ parameter, which is given as

$$Q \equiv \frac{c_s \Omega_{\text{orb}}}{\pi G \sigma},$$

in the case where we assume Keplerian motion. Here, $c_s$ denotes the sound velocity of gas, $\Omega_{\text{orb}}$ is the angular velocity around the central sink particle, and $\sigma$ is the column density of the disk. In the case that the $Q$ parameter is less than unity, a smooth disk with density perturbations becomes gravitationally unstable. In fact, the $Q$ parameter of the disk at the early phase (top) is already less than unity, so the disk is unstable (middle and bottom).

The time scale of disk instability is given by linear perturbation theory (Toomre 1964), which is given by $\Omega_{\text{orb}}^{-1} (Q^{-2} - 1)^{-1/2}$. The typical value of the $Q$ parameter in the disk at the early phase (top) is $\sim 0.5$, and the angular velocity is $\Omega_{\text{orb}} \simeq 10^{-5}$ s$^{-1}$. Thus, the perturbation growth time scale is $\sim 20$ yr, which is comparable to the time scale of the generation of the spiral structure. Thus, spiral structures seem to develop due to gravitational instability, which can be understood as the Toomre criteria. On the other hand, it takes several hundred years for another sink to be born (bottom), seemingly via the fragmentation of the spiral arms. Therefore, the formation of the sink particles in the disk cannot be understood solely by the simple $Q$ parameter argument above; nonlinear calculations are required.

3.2. Effects of Radiative Feedback

Figure 2 illustrates the edge-on view of the evolution of the central $10^4$ AU (0.05 pc) radius. Color corresponds to the fraction of H$_2$ molecules and the transparency denotes the gas density. Small spheres are the position of sink particles. Initially, the H$_2$ fraction is quite high ($y_{H_2} \sim 10^{-2}$; upper left panel). Eventually, the polar region is photodissociated as the sink particles grow (top right). Some H$_2$-rich regions remain along the equatorial plane due to self-shielding (bottom left), but these finally disappear after 0.1 Myr (bottom right).

Figure 3 illustrates the evolution of the system in the density–temperature plane. The color map indicates the frequency distribution of SPH particles on the plane. The four panels show snapshots at $t = -30$ yr, 2450 yr, 5510 yr, and 100250 yr, respectively. The distribution of SPH particles in the top left panel is similar to the well-known curve of the evolution of collapsing primordial gas in the runaway phase.
Figure 3. Four time snapshots of the density–temperature plane. The color represents the number of SPH particles dropped in the logarithmic bin on the plane. The thick solid lines show the resolution limit of this simulation, while the dashed lines represent the number density above which the SPH particles are converted into sink particles.

(e.g., Palla et al. 1983), since it corresponds to the epoch just before the first sink formation. After the first sink formation, dense gas ($\gtrsim 10^{10}$ cm$^{-3}$) is split into high temperature gas ($\lesssim 7000$ K) and low temperature gas ($\lesssim 1000$ K; top right panel). The former corresponds to the radiatively heated gas and the shock-heated gas, while the latter is the self-shielded cold gas orbiting around the sink particles. Then, the high temperature gas in the high-density region expands due to increased thermal pressure, generating a shock propagating to the low density region (bottom left panel). In fact, we also have seen the shock wave in the bottom left panel of Figure 2, marked by the white dashed curve. Finally, the dense cold gas disappears (bottom right), which means that star formation no longer proceeds.

The solid line in Figure 4 shows the time evolution of the total mass in the sink particles in the feedback run, whereas the dashed line represents that without radiative feedback effects. It is clear that the mass accretion onto sink particles is highly suppressed by radiative feedback. The total sink mass of the run with feedback at the end of the simulation ($\sim$0.1 Myr) is less than a third of that without feedback. We also find that the mass accretion rate in the feedback run is smoother than that of the no feedback run. This is because the gas in the latter case is more clumpy than in the former, due to the absence of additional heating processes provided by the ultraviolet radiation from the protostar. Figure 5 illustrates the time evolution of the mass of each sink particle. The red curves represent those in the run with feedback, while the green curves are those results without feedback. In the feedback run, we have one star more massive than $50M_\odot$ at 0.1 Myr ($\sim$57 $M_\odot$), whereas we have three stars in the range of 50–200 $M_\odot$ in the no feedback run.
the fragmentation of the disk in the no feedback case continues until much later times (∼4 × 10^4 yr) than in the feedback run (≤1500 yr), because the molecular-rich disk is not destroyed until a much later phase in the no feedback run. The difference in the number of sink particles and the minimal mass might come from such an effect. However, it is premature to draw definitive conclusions on this issue, since our results are based on only a single realization.

Figure 6 plots the sink formation time and the final mass (10^5 yr after the first sink) for all sink particles. Red plus signs denote the sinks found in feedback run, while the green crosses are the no feedback run. Basically, earlier formation leads to more massive sinks because massive sinks gather gas more efficiently than less massive ones. This trend is clear in the feedback run, while some different behaviors are found in the no feedback case. In the latter case, the gas distribution is more clumpy, which allows for sink formation at late epochs such as 20 kyr or 40 kyr after the first sink formation, as mentioned in the previous paragraph. In such cases, the growth of sinks is not affected by the sinks formed at the first episode, because the growth proceeds at positions that are relatively spatially distant from the first sinks.

3.3. Ejections

In the simulation with feedback, we find that a sink particle is kicked away from the central dense region via the gravitational N-body interaction, the so-called slingshot mechanism. Figure 7 shows the trajectories of all the sink particles within a (2 × 10^4 AU)^3 box around the central region. It is clear that one escaping sink leaves the central region (the bottom of the panel), whereas the other sinks remain within the box. Such a phenomenon has already been reported by other groups using simulations without radiative feedback effects (e.g., Smith et al. 2011). Thus, we confirm the theoretical existence of such escapers also in our numerical model with radiative feedback.

The velocity of this escaping sink is ∼4 km s^{-1} at 0.1 pc from the center of the cloud, which is marginal for escaping from a host minihalo of mass 10^6 M_⊙. The mass accretion onto this escaping particle almost stops after the ejection. Consequently, the mass of the sink is 4.4 M_⊙, much smaller than the "conventional" first stars with masses ≥100 M_⊙. It is also worth noting that the orbit of the other sinks is excited by the N-body interactions with each other, although they are not ejected. Thus, some of the sinks go through relatively low-density regions, which results in low-mass accretion rates onto these sinks.

We also remark that a star less massive than 0.8 M_⊙ is found in our higher resolution run (M_{\text{res}} = 0.1 M_⊙) with feedback, although the integrated physical time is ∼2 × 10^4 yr (Umemura et al. 2012). Considering the fact that the mass resolution and the accretion radius of the present simulations are ∼0.5 M_⊙ and 30 AU, respectively, and that other higher resolution studies with/without feedback effects report the ejection of even lower-mass stars (Umemura et al. 2012; Clark et al. 2011a, 2011b;
Susa

4. DISCUSSIONS AND CONCLUSIONS

In the presence of radiative feedback, the gas in the neighborhood of protostars is heated up significantly. This heating process occurs mainly through photodissociation of H$_2$ molecules: the formation of H$_2$ molecules works to heat the gas, since formation processes such as 3H$\rightarrow$H$_2$+H or H$^-$+H $\rightarrow$ H$_2$ + e$^-$ release latent heat. In the absence of radiative dissociation processes, collisional dissociation processes absorb thermal energy, which balances the formation heating. Thus, the net increase of H$_2$ molecules causes effective heating of the gas, while the net decrease of H$_2$ molecules results in cooling. On the other hand, the presence of strong photodissociative radiation overpowers other collisional dissociation processes. The photodissociation process does not absorb thermal energy, since the energy required to dissociate H$_2$ molecules is supplied by the radiation. Therefore, H$_2$ dissociation is no longer a cooling process. As a result, H$_2$ formation heating proceeds without hindrance in the absence of the counter process. In the present simulation, a strong L–W radiation field from the protostar is the source of this heating process in dense regions. Consequently, this heating process drives the shock wave found in Figure 2, as well as the termination of mass accretion onto the protostars.

In fact, Figure 8 illustrates the ratio between the H$_2$ formation heating rate and the adiabatic heating rate as a function of the gas temperature and density at $t = 2450$ yr. It is clear that H$_2$ formation heating is the dominant heating process at high densities ($n_H \gtrsim 10^{10}$ cm$^{-3}$, right panel). The temperature of these chemically-heated high-density regions is typically around 1000 K, and chemical heating is also important for the high-temperature region at 3000 K $\lesssim T \lesssim 7000$ K (left panel). It is also clear that the chemical heating rate is non-negligible compared with the adiabatic heating rate even at lower densities ($\sim 10^6$ cm$^{-3}$) and lower temperatures ($T \sim 300$ K); these states correspond to the dissociated polar regions. Thus, the chemical heating of H$_2$ formation plays an important role in the dynamical evolution of this system.

Heating through photoionization is also important. In fact, in similar calculations in 2D by Hosokawa et al. (2011), photoionization is the dominant heating process. These authors found the breakout of ionization fronts into the polar region and highly ionized gas heated to $> 10^4$ K. Finally, the disk is photoevaporated mainly due to the photoheating process through ionization. In the present calculation, however, ionization is not the dominant heating process. The reason simply comes from the fact that the spatial resolution of the present simulation is not enough to resolve the “initial” Strömgren sphere (Spitzer 1978). In Figure 9, we plot the ratio between the Strömgren radius $R_{ST}$ and the SPH spatial resolution $h_{SPH}$ at several densities, as a function of the protostellar mass. The three curves correspond to $n_H = 10^{10}$ cm$^{-3}$, $10^8$ cm$^{-3}$, and $10^6$ cm$^{-3}$. $R_{ST}$ is not resolved in the shaded region.
the present simulation, gas particles in the neighborhood of the source stars are always denser than \(10^{2} \text{ cm}^{-3}\), which is too high to capture the propagation of ionization fronts. Hence, only the SPH particles in the very vicinity of the source stars are heated by ionization “mildly,” \(T \lesssim 10^{4} \text{ K}\), which cannot halt the mass accretion onto the protostar and cannot cause the breakout of the D-type ionization front that is found by Hosokawa et al. (2011). In other words, this radiative feedback due to the photoheating effect is heavily underestimated in the present simulation.

As discussed in Section 2, the algorithm of mass accretion onto sink particles overestimates the mass accretion rate. Combined with the fact that feedback is underestimated, the mass accretion rate obtained in the present numerical experiment is overestimated by any account. Accordingly, the final mass of the sinks should be regarded as an upper limit on the actual mass of the first stars. If we could perform simulations at higher resolution with more realistic mass accretion conditions, the final mass could be smaller than \(\sim 50 M_\odot\). We also should keep in mind that this result comes from a numerical experiment of one realization. Thus, this upper limit should be regarded as a guide, since slightly different initial conditions could cause different results because of the chaotic nature of the system. On the other hand, at least one of the protostars cannot be less massive than \(20 M_\odot\), since radiative feedback becomes important only for \(M > 20 M_\odot\). Thus, the mass of the primary star among the first stars formed in a single minihalo will fall in the range of \(20 M_\odot\text{several } 10 M_\odot\).

We also remark that evidence for the existence of \(15–50 M_\odot\) Population III stars has been known for several years (Beers & Christlieb 2005; Frebel et al. 2005; Cayrel et al. 2004; Iwamoto et al. 2005; Lai et al. 2008; Joggerst et al. 2010; Caffau et al. 2012), consistent with the present results. In addition, however, the present results do not exclude the presence of very massive stars of \(> 130 M_\odot\) that result in pair-instability supernovae; the present simulation is a result of only one realization. Searching for evidence of strong “odd–even” effects known to be the mark of pair instability supernovae in damped Ly\(\alpha\) systems (Cooke et al. 2011) or metal poor stars (Ren et al. 2012) could provide constraints on this type of theoretical experiments.

The initial condition of the present experiment is a rigidly rotating Bonner–Ebert sphere. Although its angular momentum distribution just before the sink formation is close to that of cosmological simulations, it is not fully cosmological. In cosmological simulations, the direction of angular momentum of the disk around the primary protostar changes depending on the stages, since the gas motion is more turbulent. In addition, we need more numerical experiments starting from various initial conditions in order to obtain statistical quantities such as the initial mass function of the first stars. Thus, it is important to perform fully cosmological simulations of this sort. This work is left for future studies.

In this paper, we have investigated the suppression of mass accretion onto Population III protostars by their radiative feedback. We performed numerical experiments of the formation of the first stars using the three-dimensional radiative hydrodynamics code RSPH combined with the sink particle technique. Consequently, we find that mass accretion is suppressed significantly and that the final outcome is a multiple stellar system consisting of five stars with masses ranging from 1–60 \(M_\odot\). The fact that low-mass stars are found in this work suggests the possible existence of first stars in the local universe, although the mass of the formed stars in our simulation are still larger than \(0.8 M_\odot\).

We thank the anonymous referee for his/her careful reading and constructive comments. We also thank T. Hosokawa for providing the data of protostars and K. Omukai for a careful reading of the manuscript. This work was supported by the Ministry of Education, Science, Sports and Culture, Grant-in-Aid for Scientific Research (C), 22540295.

REFERENCES

Abel, T., Bryan, G. L., & Norman, M. L. 2000, ApJ, 540, 39
Abel, T., Bryan, G. L., & Norman, M. L. 2002, Sci, 295, 93
Bate, M. R., Bonnell, I. A., & Price, N. M. 1995, MNRAS, 277, 362
Bate, M. R., & Burkert, A. 1997, MNRAS, 288, 1060
Beers, T. C., & Christlieb, N. 2005, ARA&A, 43, 531
Bromm, V., Coppi, P. S., & Larson, R. B. 1999, ApJL, 527, L5
Bromm, V., Coppi, P. S., & Larson, R. B. 2002, ApJ, 564, 23
Caffau, E., Benoist, F., Francois, P., et al. 2012, A&A, 542, A51
Cayrel, R., Deglile, E., Spite, M., et al. 2004, A&A, 416, 1117
Clark, P. C., Glover, S. C. O., Klessen, R. S., & Bromm, V. 2011a, ApJ, 727, 110
Clark, P. C., Glover, S. C. O., Smith, R. J., et al. 2011b, Sci, 331, 1040
Cooke, R., Pettini, M., Steidel, C. C., Rudie, G. C., & Jorgenson, R. A. 2011, MNRAS, 412, 1047
Frebel, A., Aoki, W., Christlieb, N., et al. 2005, Natur, 434, 871
Fuller, T. M., & Couchman, H. M. P. 2000, ApJ, 543, 6
Galli, D., & Palla, F. 1998, A&A, 335, 403
Greif, T. H., Bromm, V., Clark, P. C., et al. 2012, MNRAS, 424, 399
Greif, T. H., Springel, V., White, S. D. M., et al. 2011, ApJ, 737, 75
Haiman, Z., Thoul, A. A., & Loeb, A. 1996, ApJ, 464, 523
Hosokawa, T., & Omukai, K. 2009, ApJ, 691, 823
Hosokawa, T., Omukai, K., Yoshida, N., & Yorke, H. W. 2011, Sci, 334, 1250
Iwamoto, N., Nomoto, K., & Maeda, K. 2005, Sci, 330, 451
Kepner, J. V., Babul, A., & Spergel, D. N. 1997, ApJ, 487, 61
Lai, D. K., Bolte, M., Johnson, J. A., et al. 2008, ApJ, 681, 1524
Martel, H., Evans, N. J., IL, & Shapiro, P. R. 2006, ApJS, 163, 122
Nakamura, F., & Umemura, M. 2001, ApJ, 548, 19
Nishi, R., & Susa, H. 1999, ApJL, 523, L103
Omukai, K. 2000, ApJ, 534, 809
Omukai, K., & Nishi, R. 1998, ApJ, 508, 141
Omukai, K., & Palla, F. 2003, ApJ, 589, 577
Palla, F., Salpeter, E. E., & Stahler, S. W. 1983, ApJ, 271, 632
Ren, J., Christlieb, N., & Zhao, G. 2012, RAA, 12, 1637
Ripamonti, E., & Abel, T. 2004, MNRAS, 348, 1019
Shapiro, P. R., & Kang, H. 1987, ApJ, 318, 32
Smith, R. J., Glover, S. C. O., Clark, P. C., Greif, T., & Klessen, R. S. 2011, MNRAS, 414, 3633
Smith, R. J., Hosokawa, T., Omukai, K., Glover, S. C. O., & Klessen, R. S. 2012, MNRAS, 424, 457
Spitzer, L. 1978, Physical Processes in the Interstellar Medium (New York: Wiley-Interscience) 333
Stacy, A., Greif, T. H., & Bromm, V. 2012, MNRAS, 422, 290
Stahler, S. W., Palla, F., & Salpeter, E. E. 1986, ApJ, 302, 590
Stancil, P. C. 1994, ApJ, 430, 360
Susa, H. 2006, PASJ, 58, 445
Susa, H., & Umemura, M. 2004, ApJ, 600, 1
Tegmark, M., Silk, J., Rees, M. J., et al. 1997, ApJ, 474, 1
Toomre, A. 1964, ApJ, 139, 1217
Turk, M. J., Abel, T., & O'Shea, B. 2009, Sci, 320, 601
Umemura, M., Susa, H., Hasegawa, K., Suwa, T., & Semelin, B. 2012, PTEP, 01A306
Wolcott-Green, J., Haiman, Z., & Bryan, G. L. 2011, MNRAS, 418, 838
Yoshida, N., Abel, T., Hernquist, L., & Sugiyama, N. 2003, ApJ, 592, 645
Yoshida, N., Omukai, K., & Hernquist, L. 2008, Sci, 321, 669
Yoshida, N., Omukai, K., Hernquist, L., & Abel, T. 2006, ApJ, 652, 6