EVOLUTION OF INTERSTELLAR CLOUDS IN LOCAL GROUP DWARF SPHEROIDAL GALAXIES IN THE CONTEXT OF THEIR STAR FORMATION HISTORIES

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ABSTRACT

We consider evolution of interstellar clouds in Local Group dwarf spheroidal galaxies (dSphs) in the context of their observed star formation histories. The Local Group dSphs generally experienced initial bursts of star formations in their formation epochs (∼ 15 Gyr ago), when hot gas originating from the supernovae can make the cold interstellar clouds evaporate. We find that the maximum size of evaporating cloud is 10 pc. Thus, clouds larger than 10 pc can survive during the initial star formation. These surviving clouds can contribute to the second star formation to produce “intermediate-age (∼ 3–10 Gyr ago) stellar populations.” Assuming that collisions between clouds induce star formation and that the timescale of the second star formation is a few Gyr, we estimate the total mass of the clouds. The total mass is about $10^4 M_\odot$, which is 1–3 orders of magnitude smaller than the typical stellar mass of a present dSph. This implies that the initial star formation is dominant over the second star formation, which is broadly consistent with the observed star formation histories. However, the variety of the dSphs in their star formation histories suggests that the effects of environments on the dSphs may be important.

Subject headings: conduction — galaxies: evolution — galaxies: ISM — galaxies: Local Group

1. INTRODUCTION

Recent observations have been revealing the properties of the Local Group dwarf spheroidal galaxies (dSphs). The dSphs have luminosities of order $10^5 - 7 L_\odot$ and are characterized by their low surface brightnesses (see Gallagher & Wyse 1994 for review).
The dSphs contain such small amounts of gas that they show no evidence of present star formation.\(^1\) Saito (1979) showed that instantaneous gas ejection from supernovae (SNe) can make the gas in proto-dSphs escape in their initial burst of star formation (see also Larson 1974). This so-called SN feedback mechanism nicely accounts for the observed scaling relations among mass, luminosity, and metallicity of each dSph (Dekel & Silk 1986; see also Hirashita, Takeuchi, & Tamura 1998).

The stellar population analyses of the Local Group dSphs show that their star formation histories are full of variety (Mateo et al. 1998; Aparicio 1999; Grebel 1999). Some dSphs has prominent intermediate-age (∼3–10 Gyr ago) stellar populations, and others have only small numbers of such populations. A dSph located close to the Galaxy tends to have poor intermediate-age stars. This may indicate that the Galaxy has affected their star formation histories through ultraviolet (UV) radiation or the Galactic wind (van den Bergh 1994).

In this paper, we aim at understanding of the evolution of interstellar medium in the Local Group dSphs in the context of their star formation histories. Though it is generally difficult to infer the physical properties of interstellar gas of the dSphs in their star formation epochs, star formation histories derived from the stellar color-magnitude diagrams (e.g., Gallagher & Wyse 1994) help us to obtain some information on the physical quantities of the gas. A merit of using the Local Group dSphs is that they are so close to the Galaxy that their star formation histories are directly inferred from their stellar populations. This work can be applied to dwarf irregular galaxies, elliptical galaxies or distant galaxies in the future.

This paper is organized as follows. First of all, in the next section, we consider the evolution of interstellar clouds in initial star formation epoch. Then, in §3, we apply the cloud-cloud collision model to star formation in the intermediate ages of dSphs. Finally, the last section is devoted to discussions.

### 2. SURVIVAL OF CLOUDS IN PROTO-DWARF GALAXIES

The collapse of gas in a proto-dSph induces the initial burst of star formation. This initial burst occurs in the dynamical timescale determined by the dark matter potential (∼10⁷ yr). During the burst, the hot gas (temperature of $T \sim 10^6$ K) originating mainly from SNe (McKee & Ostriker 1977) contributes to heating of interstellar gas through

\(^1\)In fact, a dwarf galaxy with present star formation is not generally called dSph.
thermal conduction (e.g., Draine & Giuliani 1984). In this section, we examine the effect of the thermal conduction.

2.1. Evaporation Timescale of Clouds

It is widely accepted that interstellar medium is a cloudy fluid (e.g., Elmegreen 1991). Interstellar medium is multiphase gas with various temperature and number density (McKee & Ostriker 1977). Here, we simply consider two-phase interstellar gas composed of hot \( T \sim 10^6 \) K diffuse gas and cool \( T < 10^4 \) K clouds. For the evolution of interstellar medium in the context of multiphase interstellar medium, see e.g., Fujita, Fukumoto, & Okoshi (1996).

The hot gas originating from the successive SNe can heat cool interstellar clouds. The cool gas may finally evaporate. The evaporating gas as well as the hot gas escapes freely out of the proto-dwarf galaxy, since the thermal energy at \( T \sim 10^6 \) K is much larger than the gravitational potential. The timescale for the gas to escape out of the galaxy is estimated by crossing time \( t_{\text{cross}} \) defined by

\[
t_{\text{cross}} \equiv \frac{R_{\text{dSph}}}{c_s},
\]

where \( c_s \) is the sound speed of the hot gas (typically 100 km s\(^{-1}\)) and \( R_{\text{dSph}} \) is the typical size of the proto-dSph. Thus, the crossing time is estimated by

\[
t_{\text{cross}} \simeq 1.0 \times 10^7 \left( \frac{R_{\text{dSph}}}{1 \text{ kpc}} \right) \left( \frac{c_s}{100 \text{ km s}^{-1}} \right)^{-1} \text{ yr}.
\]

Hirashita, Takeuchi, & Tamura (1998) also estimated the crossing time. The estimation from the view point of thermal energy also provides the timescale of \( 10^6-7 \) yr for the escape of hot gas from dwarf galaxies (Yoshii & Arimoto 1987; Nath & Chiba 1995).

Now we discuss the process of evaporation. After treating the conservation of mass and energy, Cowie & McKee (1977) derived the typical mass loss rate \( \dot{m} \) of a cool cloud embedded in the hot medium as

\[
\dot{m} = \frac{16\pi \mu m_H \kappa_h R_c}{25k_B},
\]

where \( \mu \) is the mean weight of a particle normalized by the mass of a hydrogen atom \( (m_H) \), \( R_c \) is the radius of the cloud [a spherical cool \( T \ll T_h \), where \( T_h \) is the temperature of the hot gas) cloud is assumed], \( k_B \) is the Boltzmann constant, and \( \kappa_h \) is the thermal
conductivity estimated at the temperature of \( T_h \) and the electron number density of \( n_h \) (the electron number density of the hot gas). The thermal conductivity is expressed as

\[
\kappa_h = 1.8 \times 10^{-5} \frac{T_h^{5/2}}{\ln \Lambda} \text{ erg s}^{-1} \text{ deg}^{-1} \text{ cm}^{-1},
\]

where \( \ln \Lambda \) is the Coulomb logarithm which is a function of electron number density and electron temperature (Spitzer 1956; Cowie & McKee 1977). Using \( \dot{m} \), the timescale for the evaporation, \( t_{\text{evap}} \), of the cloud is estimated as

\[
t_{\text{evap}} = \frac{m_c}{\dot{m}} = 1.4 \times 10^7 \left( \frac{\bar{n}_c}{1 \text{ cm}^{-3}} \right) \left( \frac{R_c}{10 \text{ pc}} \right)^2 \left( \frac{T_h}{10^6 \text{ K}} \right)^{-5/2} \left( \frac{\ln \Lambda}{30} \right) \text{ yr},
\]

where \( \bar{n}_c \) is the mean number density of gas in the cloud and \( m_c \) is the mass of the cloud estimated by

\[
m_c = \frac{4\pi}{3} R_c^3 \mu m_H \bar{n}_c.
\]

According to Larson (1974), hot gas fills more than half of the dwarf galaxy in a short timescale (\(< 10^6 \) yr) if successive and multiple SNe are considered. Thus, our picture of continuous evaporation during the crossing time is justified. However, we should note that the timescale is largely dependent on stellar initial mass function.

### 2.2. Cooling Timescale of the Hot Gas

In this subsection, we estimate the cooling timescale \( t_{\text{cool}} \). The cooling timescale is expressed by

\[
t_{\text{cool}} = \frac{3k_B T_h}{2n_h \Lambda_{\text{cool}}(T_h)},
\]

where \( n_h \) is the number density of electrons in the hot gas and \( \Lambda_{\text{cool}}(T_h) \) is the cooling function as a function of temperature. The cooling function is composed of two contributions as

\[
\Lambda_{\text{cool}}(T_h) = \Lambda_{\text{ff}}(T_h) + \Lambda_{\text{line}}(T_h),
\]

where \( \Lambda_{\text{ff}} \) and \( \Lambda_{\text{line}} \) represent the cooling rates through free-free radiation and through metal-line emission, respectively. According to Gaetz & Salpeter (1983), the metal cooling is estimated as

\[
\Lambda_{\text{line}}(T_h = 10^6 \text{ K}) \simeq 1.3 \times 10^{-22} \zeta \text{ erg s}^{-1} \text{ cm}^3,
\]
where $\zeta$ is the metallicity normalized by the solar system abundance (see also Raymond, Cox, & Smith 1976 for the cooling function). On the other hand, the free-free cooling function is estimated as

$$\Lambda_{\text{ff}}(T_h) \simeq 2 \times 10^{-24} \left( \frac{T_h}{10^6 \text{ K}} \right)^{1/2} \text{erg s}^{-1} \text{cm}^3 \quad (10)$$

(Rybicki & Lightman 1979).

If we assume $\zeta \sim 0.01$, which is a typical value for stellar metallicity observed in the present dSphs (Aaronson & Mould 1985; Buonanno et al. 1985), $\Lambda_{\text{line}}(T_h = 10^6 \text{ K}) \simeq 1.3 \times 10^{-24} \text{erg s}^{-1} \text{cm}^3$. The resulting cooling timescale becomes

$$t_{\text{cool}} \simeq 2 \left( \frac{T_h}{10^6 \text{ K}} \right) \left( \frac{n_h}{10^{-3} \text{ cm}^{-3}} \right)^{-1} \left( \frac{\Lambda_{\text{cool}}}{3 \times 10^{-24} \text{erg s}^{-1} \text{cm}^{-1}} \right)^{-1} \text{Gyr.} \quad (11)$$

Hence, the cooling timescale is much longer than the crossing timescale (eq. 2). This means that the effect of cooling of the hot gas can be neglected. The present stellar metallicity of the dSph sample is at most $\zeta \sim 0.1$ if we consider the error range. Adopting this upper limit value, we obtain the shorter value of the cooling time of $\sim 0.5 \text{ Gyr}$. Even in this case, the cooling time is much shorter than the crossing time.

We have considered uniform hot gas distribution. The confinement of the hot gas might be possible. With the confinement, the density of the hot gas can be so high that the cooling timescale becomes shorter than the crossing timescale. However, since the crossing time of the confined region also becomes shorter than that in the previous estimation, it seems difficult to realize the shorter cooling timescale than the crossing time. Once the hot gas crosses a dense region, the hot gas blows away easily. Thus, it is reasonable to assume the longer cooling time than the crossing time. We note that if a physically reasonable confining mechanism of the hot gas is found the estimation in this section may need modification.

The mixing of the hot gas with the warm gas can make the cooling timescale short. The mixing produces the gas with temperature of $\sim 10^5 \text{ K}$ (Begelman & Fabian 1990; Slavin, Shull, & Begelman 1993), at which the line cooling rate becomes an order of magnitude larger than that at $\sim 10^6 \text{ K}$ (Gaetz & Salpeter 1983). Thus, the cooling timescale may be shorter by an order of magnitude than the previous estimate. However, even in this case, the cooling time is larger than the crossing time.

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2We note that we should consider the metallicity of gas, not stars. Though the lack of gas content in dSphs makes it impossible to know the metallicity of their gas, we imagine that the gas metallicity did not exceed 0.1 from the present gas content in low-mass dwarf irregular galaxies.
2.3. Condition for Survival of a Cloud

Here, we estimate the size of a cloud that survives the evaporation. Since the evaporation process is effective in the timescale of $t_{\text{cross}}$, the condition for the survival is expressed by

$$t_{\text{evap}} > t_{\text{cross}}.$$  \hfill (12)

From equations (2) and (5), the above condition is written as

$$R_{c} > R_{\text{crit}} \equiv 8 \left(\frac{\bar{n}_{c}}{1 \text{ cm}^{-3}}\right)^{-1/2} \left(\frac{T_{h}}{10^6 \text{ K}}\right)^{5/4} \left(\frac{\ln \Lambda}{30}\right)^{-1/2} \times \left(\frac{R_{\text{dSph}}}{1 \text{ kpc}}\right)^{1/2} \left(\frac{c_{s}}{100 \text{ km s}^{-1}}\right)^{-1/2} \text{ pc},$$  \hfill (13)

where we define the critical radius for the survival, $R_{\text{crit}}$. Thus, the cloud larger than $R_{\text{crit}} \sim 10$ pc can survive the thermal conduction during the initial burst of star formation.

The motions of interstellar clouds produce velocity shear between the clouds and ambient hot gas. The shear may lead to the Kelvin-Helmholtz (K-H) instability (Chandrasekhar 1961). Since the growth timescale of the K-H instability is as short as the conduction timescale (Appendix A) for $R_{c} \sim 10$ pc (for the notation in Appendix A, $R_{c} = \lambda$), the K-H instability as well as the conduction can determine the minimum mass of the clouds.

3. CLOUD-CLOUD COLLISION

One of the direct tests of survival of the cloud in the initial star formation in proto-dSphs is to examine star formation histories of the dSphs. Their star formation histories are investigated from the stellar population analyses (Gallagher & Wyse 1994; Mateo 1998). The second star formation produced what is called “intermediate-age stellar populations” (e.g., Gallagher & Wyse 1994). There are evidences of second star formation in each Local Group dSphs. Assuming that the second star formation is due to the surviving clouds within a dSph, we examine the physical properties of the clouds.

Before the examination, we should fix the mechanism of the second star formation. We assume that star formation is induced by collisions between clouds. This physically means that the compression of clouds during the collision makes free-fall time and cooling time of
the clouds shorter (the physical process is described in Kumai, Basu, & Fujimoto 1993), which leads to formation of dense molecular clouds and finally to star formation. Even if a cloud has formed stars, the shell of a stellar wind or a supernova remnant associating with the cloud induce star formation of another cloud which collides with it (see also Roy & Kunth 1995). The idea that cloud-cloud collisions induce star formation has a long history (e.g., Field & Saslaw 1965). We note that another possible mechanism of star formation should be investigated in future works.

The cloud-cloud collision timescale $t_{\text{coll}}$ can be estimated by

$$t_{\text{coll}} \simeq \frac{1}{N\sigma V},$$

(14)

where $N$ is the number of clouds per unit volume, $\sigma$ is the geometrical cross section of a cloud, expressed as $\sigma = \pi R_c^2$, and $V$ is the velocity of a cloud. In fact, $N\sigma V$ should be written as $\langle N\sigma V \rangle$, where $\langle \cdot \rangle$ means that the physical quantity is averaged for all the clouds. The number of the clouds is the largest in the region where the gravitational potential well is the deepest. The size of the deepest region is typically estimated by the core radius (typically 100–1000 pc for dSphs; e.g., Mateo 1998). Thus, the physical quantities in this section represents the typical values within the core radius.

The velocity $V$ is determined by the virial-equilibrium value in gravitational potential of a dSph ($\sim 10$ km s$^{-1}$). If we give a typical collision timescale, we can estimate the number density of clouds in a dSph by

$$N \simeq \frac{1}{\pi R_c^2 V t_{\text{coll}}}. $$

(15)

Observationally, the duration of second star formation seems a few Gyr (van den Bergh 1994; Mateo et al. 1998; Grebel 1999). Based on the assumption that the timescale of star formation is determined by the collision timescale of interstellar clouds, $t_{\text{coll}}$ should be $\sim$ a few Gyr. Thus, the following estimation for $N$ is possible according to equation (15):

$$N \simeq 1.0 \times 10^{-7} \left( \frac{R_c}{10 \text{ pc}} \right)^{-2} \left( \frac{V}{10 \text{ km s}^{-1}} \right)^{-1} \left( \frac{t_{\text{coll}}}{3 \text{ Gyr}} \right)^{-1} \text{ pc}^{-3},$$

(16)

where we estimated the size of the cloud by $R_{\text{crit}}$ estimated in §2.3. This value of $N$ corresponds to the mean cloud-cloud interval of $\sim 200$ pc. We note that $R_{\text{crit}}$ is the lower limit of the size of a cloud which survived the first star formation. If the typical size is larger, $N$ becomes smaller and the typical interval between clouds gets larger.

This argument about the cloud size becomes clear if we introduce the mass spectrum of clouds. The mass spectrum $\mathcal{N}(m)dm$ is defined by the number density of clouds of
mass between \( m \) and \( m + dm \). If we assume that the spectrum can be expressed by the power-law, \( N(m) \propto m^{-p} \), and that mass \( (m) \) and size \( (R) \) of any cloud is related by \( m \propto R^3 \) (the constant mass density of any cloud), we obtain the size distribution of the clouds: 
\[ N(R)dR \propto R^{-3p+2}dR. \]
Using this form, we calculate the mean size of the clouds, \( \langle R \rangle \), as
\[
\langle R \rangle = \frac{\int_{R_{\text{crit}}}^{R_{\text{up}}} R N(R)dR}{\int_{R_{\text{crit}}}^{R_{\text{up}}} N(R)dR}
= \left( \frac{3p - 4}{3p - 3} \right) \left[ \frac{(R_{\text{up}}/R_{\text{crit}})^{-3p+4} - 1}{(R_{\text{up}}/R_{\text{crit}})^{-3p+3} - 1} \right] R_{\text{crit}},
\]
where \( R_{\text{up}} \) is the upper cutoff of the size. Setting as \( R_{\text{up}} \gg R_{\text{crit}} \), we obtain \( \langle R \rangle \sim R_{\text{crit}} \) for \( p > 1 \). In this case, estimation of \( R_c \) with \( R_{\text{crit}} \) (eq. [16]) is justified. Field & Saslaw (1965) suggested \( p = 3/2 \) to explain observed star formation rate.

We further estimate the total mass of surviving clouds. The total mass \( M_c \) can be estimated by
\[
M_c \approx \frac{4\pi}{3} R_{\text{dSph}}^3 N m_c
\approx 5 \times 10^4 \left( \frac{R_{\text{dSph}}}{1 \text{ kpc}} \right)^3 \left( \frac{\bar{n}_c}{1 \text{ cm}^{-3}} \right) \left( \frac{R_c}{10 \text{ pc}} \right) \left( \frac{V}{10 \text{ km s}^{-1}} \right)^{-1} \left( \frac{t_{\text{coll}}}{3 \text{ Gyr}} \right)^{-1} M_{\odot},
\]
where \( m_c \) is defined in equation (6). Thus, the total mass of the stellar population, which is formed in the second star formation epoch (so-called intermediate age), is roughly \( \sim 10^4 M_{\odot} \) (or less, since we can hardly expect all the gas is converted to stellar mass). This is 1–3 orders of magnitude smaller than the typical stellar mass of the dSphs (\( \sim 10^5–7 M_{\odot} \)). Indeed, the number of the intermediate-age population is much smaller than the old stellar population for the dSph, though the Carina dSph has prominent intermediate-age populations. (e.g., Mateo 1998). We will comment on the exception later in §4.

Here, we should mention the effect of SNe during the second star formation. As shown above, the second star formation is not so active as the initial burst of star formation. Thus, once cloud size is determined during the initial burst through the thermal conduction, the SN heating in the intermediate age has little influence on the clouds. Thus, we can reasonably ignore the SN heating in the intermediate age.

4. DISCUSSIONS AND IMPLICATIONS

We have inferred the evolution of interstellar clouds in the Local Group dSphs from their observed star formation histories. Owing to the hot gas supplied by initial star
formation, small interstellar clouds evaporates. However, clouds larger than \( \sim 10 \) pc can survive during the burst of star formation. The surviving clouds contribute to second star formation to form so called “intermediate-age stellar populations.”

There are observational evidences that the second star formation occurred in the intermediate age (\( \sim 3–10 \) Gyr ago). The timescale of the second star formation is typically a few Gyr (e.g., Mateo et al. 1998). Assuming that star formations are induced by cloud-cloud collisions, the collision timescale should be a few Gyr to realize the observed timescale of the second star formation. Since the collision timescale is related to the number of clouds, the number is constrained. The expected number density of clouds is typically \( 1.0 \times 10^{-7} \) pc\(^{-3} \), which indicates that the total mass of gas contributing to the second star formation is typically \( 10^4 M_\odot \). This is 1–3 orders of magnitude smaller than the observed stellar mass of the Local Group dSphs. This indicates that almost all the stars in the dSphs are formed in the initial star formation.

Recently, Hirashita, Takeuchi, & Tamura (1998) suggested that the luminous mass of a dSph is determined by the depth of dark matter potential. Their suggestion is true if the first star formation is dominant in the star formation histories of dSphs. Indeed, our cloud-cloud collision model implies that the second stellar population is not dominant in mass.

We should also consider environmental effects. Van den Bergh (1994) suggested that environmental effects on dSphs may be important for their star formation histories. He showed that the star formation histories of the Local Group dwarf galaxies correlate with the Galactocentric distances (the distances from the Galaxy): Dwarf galaxies near the Galaxies, such as Ursa Minor and Draco contain only a little fraction of intermediate-age or recent stellar population, while there is observational evidences of recent star formations in distant dwarf galaxies. In the same paper, he also suggested that star formations in the dwarf galaxies are affected by the existence of UV radiation or the wind from the Galaxy. Hirashita, Kamaya, & Mineshige (1997) showed that the Galactic wind can strip the gas of nearby dwarf galaxies. In the epoch of initial burst, the effects of OB star radiations and SNe in dSphs are much stronger than environmental effects and determine the structure of dSphs (Hirashita, Takeuchi, & Tamura 1998). However, after the first star formation, such effects become weak, so that the dominant factor to determine the physical condition is environmental effects such as the Galactic wind (Hirashita, Kamaya, & Mineshige 1997). Thus, the environmental effect may be responsible for the physical nature of the second star formation. Since Ursa Minor and Draco are located closer to the Galaxy than other dSphs, they are easily affected by the ram pressure of Galactic wind which strips the gas of them. Therefore, the second star formation is less prominent in these two dSphs than other dSphs.
(van den Bergh 1994). Recently, the star formation histories of the companion dSphs of M31 has begun to be made clear (e.g., Armandroff, Davies, & Jacoby 1998). The increase of the number of sample dSphs will contribute to test of the environmental effects. Here we should note that Einasto, Saar, & Kaasik (1974) pointed out environmental effects on structures of galaxies.

Another environmental effect is possible. The burst of star formation may be induced by infall of intergalactic gas clouds. If a cloud is captured in a gravitational potential well of a dSph, the cloud may form stars. Hirashita, Kamaya, & Mineshige (1997) pointed out that the present activity of star formation in the Magellanic Clouds may be due to such infall of gas. However, since potential wells of dSphs are much shallower than those of the Magellanic Clouds, it seems difficult for dSphs to capture the intergalactic clouds.

Contrary to most dSphs, the Carina dSph shows a burst of star formation in the intermediate age. One possibility to explain this peculiar nature of the galaxy is UV radiation field. It is suggested that UV radiation field suppresses formation of dwarf galaxies (Babul & Rees 1992; Efstathiou 1992). The UV from the Galaxy may have suppressed the formation of Carina. When the UV field at the galaxy became weak in the intermediate age, Carina experienced a burst of star formation. To show that this is true, we also explain why most of dSphs did not suffer the suppression from UV radiation. The future work on the history of UV radiation field in the Local Group may provide us a hint to solve this problem.

We note that the cloud-cloud collision model is applicable to the star formation histories of dwarf irregular galaxies, giant elliptical galaxies or distant galaxies. For the application of the collision model to dense molecular clouds in a high-redshift object, see e.g., Ohta et al. (1998).

The following points remain to be solved in this paper:

[1] What determines the number of stars formed in the initial burst of star formation? This is the problem raised also in Hirashita, Takeuchi, & Tamura (1998).

[2] What determines the number of clouds that contribute to the intermediate-age stellar populations? This question is related to the mass spectrum of interstellar clouds (§3).

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APPENDIX

A. THE KELVIN-HELMHOLTZ INSTABILITY

The Kelvin-Helmholtz (K-H) instability has been discussed in various astrophysical contexts. For example, Miyahata & Ikeuchi (1995) discussed the stability of protogalactic cloud against the K-H instability (see also Murray et al. 1993). The K-H instability is also applied to interstellar physics: Klein, McKee, & Collella (1994) examined the interstellar clouds, while Fleck (1984) and Kamaya (1996) investigated molecular clouds. All of these works assume that dense clouds are embedded in a diffuse medium. Since the dense medium generally moves at the velocity determined by the depth of gravitational potential, there is generally a relative motion between the dense and diffuse media. The presence of the relative motion makes it possible to discuss the K-H instability (e.g., Chandrasekhar 1961).

In §2, we considers interstellar clouds embedded in hot tenuous gas originating from successive SNe. Since clouds in a galaxy generally move at the velocity determined by the gravitational potential of the galaxy, it is necessary to examine the timescale of the growth of the K-H instability. The growth rate (Ω) in the linear regime at a flat interface between the cloud and ambient hot gas can be expressed as

\[ \omega = k \left( \frac{\rho_c \rho_h}{\rho_c + \rho_h} \right)^{1/2} U, \]  

where \( k \) is a wavenumber of a mode (\( k = \frac{2\pi}{\lambda} \), where \( \lambda \) is the wavelength), \( \rho_c \) and \( \rho_h \) are, respectively, the mass densities of cloud and hot gas, and \( U \) is the relative velocity (Drazin & Reid 1981). Since \( \rho_c \gg \rho_h \), we obtain the following estimation for a typical proto-dwarf galaxy considered in §2:

\[ \omega \simeq 6.4 \times 10^{-14} \left( \frac{\lambda}{10 \text{ pc}} \right)^{-1} \left( \frac{\rho_h/\rho_c}{10^{-3}} \right)^{1/2} \left( \frac{U}{10 \text{ km s}^{-1}} \right) \text{s}^{-1}. \]  

The timescale of the growth, \( t_{KH} \), is estimated by

\[ t_{KH} \equiv \frac{2\pi}{\omega} \simeq 3.1 \times 10^7 \left( \frac{\lambda}{10 \text{ pc}} \right) \left( \frac{\rho_h/\rho_c}{10^{-3}} \right)^{-1/2} \left( \frac{U}{10 \text{ km s}^{-1}} \right)^{-1} \text{yr}, \]  

which is comparable with the evaporation time defined in §2. Thus, the instability may determine the size of clouds (see e.g., Kamaya 1998 for stabilizing effects).
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