ISM Dynamics and Star Formation

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many thanks to

- Robi Banerjee (ITA)
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Agenda

- phenomenology
  - what we need to explain
- dynamic star formation theory
  - gravity vs. turbulence (and all the rest)
- examples and predictions
  - formation of molecular clouds in galactic disks (H₂ & CO chemistry)
  - universal IMF: importance of turbulence and thermodynamics
phenomenology
correlation between $H_2$ and $H_1$

Compare $H_2$ - $H_1$ in M33:

- $H_2$: BIMA-SONG Survey, see Blitz et al.
- $H_1$: Observations with Westerbork Radio T.

$H_2$ clouds are seen in regions of high $H_1$ density (in spiral arms and filaments)

(Deul & van der Hulst 1987, Blitz et al. 2004)
Star formation in Orion

We see

- **Stars** (in visible light)
- Atomic hydrogen (in $\text{H}_\alpha$ -- red)
- Molecular hydrogen $\text{H}_2$ (radio emission -- color coded)
The Orion molecular cloud is the birthplace of several young embedded star clusters. The Trapezium cluster is only visible in the IR and contains about 2000 newly born stars.
stars form in clusters
stars form in molecular clouds
(proto)stellar feedback is important

Trapezium Cluster (detail)

(color composite J,H,K by M. McCaughrean, VLT, Paranal, Chile)
Trapezium Cluster: Central Region

Ionizing radiation from central star \( \Theta 1C \text{ Orionis} \)

**Proplyds:** Evaporating ``protoplanetary'' disks around young low-mass protostars

(images: Doug Johnstone et al.)
stellar mass function

- stars seem to follow a universal mass function at birth --> IMF

Orion, NGC 3603, 30 Doradus
(Zinnecker & Yorke 2007)

(Kroupa 2002)
nearby molecular clouds

scales to same scale

(Orion - Perseus - Ophiuchus - Taurus - Pipe)

10 pc near molecular clouds

(from A. Goodman)
nearby molecular clouds

scales to same scale (from A. Goodman)

study more closely
LOS Geschwindigkeitsverteilung in Perseus

velocity cube from Alyssa Goodman: COMPLETE survey
what we need to consider ...

*correlation* between large and small scales in galaxy (stars “know” where to form and when)

all stars form in *molecular cloud* complexes (star formation linked to molecular cloud formation)

molecular clouds are *turbulent* (understand turbulence to understand star formation)

stars form in *clusters* (importance of dynamical interactions during formation)

star formation has *universal* characteristics (e.g. initial mass function)
basic idea
dynamical SF in a nutshell

- interstellar gas is highly *inhomogeneous*
  - *gravitational instability*
  - *thermal instability*
  - *turbulent compression* (in shocks $\delta \rho/\rho \propto M^2$; in atomic gas: $M \approx 1...3$)

- cold *molecular clouds* can form rapidly in high-density regions at **stagnation points of convergent large-scale flows**
  - chemical *phase transition*: atomic $\rightarrow$ molecular
  - process is *modulated* by large-scale *dynamics* in the galaxy

- inside *cold clouds*: turbulence is highly supersonic ($M \approx 1...20$)
  $\rightarrow$ *turbulence* creates large density contrast,
  *gravity* selects for collapse

  **GRAVOTUBULENT FRAGMENTATION**

- *turbulent cascade*: local compression *within* a cloud provokes collapse
  $\rightarrow$ formation of individual *stars* and *star clusters*

(e.g. Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)
Turbulent cascade

Inertial range: **scale-free behavior of turbulence**

„Size“ of inertial range:

\[ \frac{L}{\eta_K} \approx \text{Re}^{3/4} \]

Kolmogorov (1941) theory incompressible turbulence
Turbulent cascade

Shock-dominated turbulence

\[ \log E \]

\[ \log k \]

energy input scale

energy dissipation scale

\[ L^{-1} \]

\[ \eta_K^{-1} \]

scale-free behavior of turbulence

inertial range: 

\[ \frac{L}{\eta_K} \approx \text{Re}^{3/4} \]

„size“ of inertial range:
Turbulent cascade in ISM

Molecular clouds: \( \sigma_{\text{rms}} \approx \text{several km/s} \), \( M_{\text{rms}} > 10 \), \( L > 10 \text{ pc} \)

Massive cloud cores: \( \sigma_{\text{rms}} \approx \text{few km/s} \), \( M_{\text{rms}} \approx 5 \), \( L \approx 1 \text{ pc} \)

Dense protostellar cores: \( \sigma_{\text{rms}} << 1 \text{ km/s} \), \( M_{\text{rms}} \leq 1 \), \( L \approx 0.1 \text{ pc} \)

Energy source & scale: \( \eta_{K}^{-1} \), \( L^{-1} \)

Sonic scale

Supersonic

Subsonic

Energy source & scale NOT known (supernovae, winds, spiral density waves?)

Dissipation scale not known (ambipolar diffusion, molecular diffusion?)
Density structure of MC’s

(Motte, André, & Neri 1998)

Molecular clouds are highly inhomogeneous.

Stars form in the densest and coldest parts of the cloud.

ρ-Ophiuchus cloud seen in dust emission.

Let’s focus on a cloud core like this one.
Evolution of cloud cores

- How does this core evolve? Does it form one single massive star or cluster with mass distribution?

- Turbulent cascade "goes through" cloud core
  --> NO scale separation possible
  --> NO effective sound speed

- Turbulence is supersonic!
  --> produces strong density contrasts: \( \frac{\delta \rho}{\rho} \approx M^2 \)
  --> with typical \( M \approx 10 \) --> \( \frac{\delta \rho}{\rho} \approx 100 \)

- many of the shock-generated fluctuations are Jeans unstable and go into collapse

- --> expectation: core breaks up and forms a cluster of stars
Evolution of cloud cores

indeed $\rho$-Oph B1/2 contains several cores ("starless" cores are denoted by $\times$, cores with embedded protostars by $\star$)

(Motte, André, & Neri 1998)
What happens to distribution of cloud cores?

Two extreme cases:

1. turbulence dominates energy budget: 
   \[ \alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} > 1 \]
   --> individual cores do not interact
   --> collapse of individual cores dominates stellar mass growth
   --> loose cluster of low-mass stars

2. turbulence decays, i.e. gravity dominates: 
   \[ \alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} < 1 \]
   --> global contraction
   --> core do interact while collapsing
   --> competition influences mass growth
   --> dense cluster with high-mass stars
turbulence creates a hierarchy of clumps
as turbulence decays locally, contraction sets in
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While region contracts, individual clumps collapse to form stars.
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individual clumps collapse to form stars
individual clumps collapse to form stars
in dense clusters, clumps may merge while collapsing
--> then contain multiple protostars

$\alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} < 1$
in **dense clusters**, clumps may merge while collapsing
-->
then contain multiple protostars
in *dense clusters*, clumps may merge while collapsing
--> then contain multiple protostars
in *dense clusters*, competitive mass growth becomes important
in *dense clusters*, competitive mass growth becomes important
in dense clusters, \( N \)-body effects influence mass growth
low-mass objects may become ejected --> accretion stops
feedback terminates star formation
result: *star cluster*, possibly with HII region
result: *star cluster* with H\textsc{ii} region
applications
two examples

- formation of molecular clouds in the disk of the Milky Way
  - timescales
  - dynamic properties
  - x-factor
- formation of star clusters inside these clouds
  - IMF
molecular cloud formation
molecular cloud formation

- star formation on galactic scales
  → missing link so far:
    formation of molecular clouds

- questions
  - where and when do molecular clouds form?
  - what are their properties?
  - how does that correlation to star formation?
  - global correlations? → Schmidt law
Molecular clouds form at stagnation points of large-scale convergent flows, mostly triggered by global (or external) perturbations.

(Deul & van der Hulst 1987, Blitz et al. 2004)
modeling galactic SF

SPH calculations of self-gravitating disks of stars and (isothermal) gas in dark-matter potential, sink particles measure local collapse --> star formation

We find correlation between star formation rate and gas surface density:

$$\Sigma_{SFR} \propto \Sigma_{\text{gas}}^{1.5}$$

observed Schmidt law

\[ \Sigma_{\text{SFR}} \propto \Sigma_{\text{gas}}^{1.5} \]

in both cases:

(from Kennicutt 1998)
correlation with large-scale perturbations

density/temperature fluctuations in warm atomar ISM are caused by thermal/gravitational instability and/or supersonic turbulence.

some fluctuations are dense enough to form $H_2$ within “reasonable time” → molecular cloud

(Glover & Mac Low 2007a,b)

e external perturbuations (i.e. potential changes) increase likelihood
molecular cloud formation

(from Dobbs, Glover, Clark, Klessen 2008)
molecular cloud formation

(Dobbs & Bonnell 2007)
molecular cloud formation

molecular gas fraction as function of time

molecular gas fraction as function of density

(Dobbs et al. 2008)
molecular cloud formation

molecular gas fraction of fluid element as function of time

molecular gas fraction as function of density

(Dobbs et al. 2008)
observed timescales

Tamburro et al. (2008)

Fig. 1.— NGC 5194: the 24 μm band image is plotted in color scale; the H I emission map is overlayed with green contours.
observed timescales

Fig. 5.— Histogram of the time scales $t_{\text{HI} \to 24 \mu m}$ derived from the fits in Figure 4 and listed in Table. 2 for the 14 sample galaxies listed in Table. 1. The timescales range between 1 and 4 Myr for almost all galaxies.

Tamburro et al. (2008)
calculated timescales

Dobbs et al. (2008)

Figure 16. This histogram gives the distribution of timescales over which the gas reaches certain molecular gas fractions. The timescales denote the time for the H$_2$ fraction of a particle to increase from 0.001 to 0.01, 0.01 to 0.1 and 0.1 to 0.5, as indicated.
molecular cloud formation
zooming in ...
image from Alyssa Goodman: COMPLETE survey
(movie from Christoph Federrath)
Large-eddy simulations

- We use **LES** to model the large-scale dynamics
- Principal problem: only large scale flow properties
  - Reynolds number: \( \text{Re} = \frac{LV}{\nu} \) \( (\text{Re}_{\text{nature}} \gg \text{Re}_{\text{model}}) \)
  - dynamic range much smaller than true physical one
- need **subgrid model** (in our case simple: only dissipation)
- but what to do for more complex when processes on subgrid scale determine large-scale dynamics
  (chemical reactions, nuclear burning, etc)
- Turbulence is “space filling” --> difficulty for AMR (don’t know what criterion to use for refinement)
- How **large** a Reynolds number do we need to catch basic dynamics right?
experimental set-up

- AMR MHD \((B = 2 \mu G)\)
- stochastic forcing
  (Ornstein-Uhlenbeck)
- self-gravity
- time-dependent chemistry
- cooling & heating processes
  --> thermodynamics done right!
- gives you mathematically well defined boundary conditions
  --> good for statistical studies
chemical model 0

- 32 chemical species
  - 17 in instantaneous equilibrium:
    
    \[ \text{H}^-, \text{H}_2^+, \text{H}_3^+, \text{CH}^+, \text{CH}_2^+, \text{OH}^+, \text{H}_2\text{O}^+, \text{H}_3\text{O}^+, \text{CO}^+, \text{HOC}^+, \text{O}^-, \text{C}^- \text{ and } \text{O}_2^- \]
  
  - 19 full non-equilibrium evolution
    
    \[ \text{e}^-, \text{H}^+, \text{H}, \text{H}_2, \text{He}, \text{He}^+, \text{C}, \text{C}^+, \text{O}, \text{O}^+, \text{OH}, \text{H}_2\text{O}, \text{CO}, \text{C}_2, \text{O}_2, \text{HCO}^+, \text{CH}, \text{CH}_2 \text{ and } \text{CH}_3^+ \]

- 218 reactions
- various heating and cooling processes

(Glover, Federrath, Mac Low, Klessen, 2010)
<table>
<thead>
<tr>
<th>Process</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>Cooling:</strong></td>
<td></td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H) – Abrahamsson, Krems &amp; Dalgarno (2007)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H₂) – Schroder et al. (1991)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (e⁻) – Johnson et al. (1987)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H⁺) – Roueff &amp; Le Bourlot (1990)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H₂) – Flower &amp; Launay (1977)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H, T &gt; 2000 K) – Keenan et al. (1986)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (e⁻) – Wilson &amp; Bell (2002)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H) – Abrahamsson, Krems &amp; Dalgarno (2007)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H₂) – see Glover &amp; Japppen (2007)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (e⁻) – Bell, Berrington &amp; Thomas (1998)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates (H⁺) – Pequignot (1990, 1995)</td>
</tr>
<tr>
<td>H₂ rovibrational lines</td>
<td>Le Bourlot, Pineau des Forêts &amp; Flower (1999)</td>
</tr>
<tr>
<td>CO and H₂O rovibrational lines</td>
<td>Neufeld &amp; Kaufman (1993); Neufeld, Lepp &amp; Melnick (1995)</td>
</tr>
<tr>
<td>OH rotational lines</td>
<td>Pavlovski et al. (2002)</td>
</tr>
<tr>
<td>Gas-grain energy transfer</td>
<td>Hollenbach &amp; McKee (1989)</td>
</tr>
<tr>
<td>Recombination on grains</td>
<td>Wolfire et al. (2003)</td>
</tr>
<tr>
<td>Atomic resonance lines</td>
<td>Sutherland &amp; Dopita (1993)</td>
</tr>
<tr>
<td>H collisional ionization</td>
<td>Abel et al. (1997)</td>
</tr>
<tr>
<td>H₂ collisional dissociation</td>
<td>See Table B1</td>
</tr>
<tr>
<td>Compton cooling</td>
<td>Cen (1992)</td>
</tr>
<tr>
<td><strong>Heating:</strong></td>
<td></td>
</tr>
<tr>
<td>Photoelectric effect</td>
<td>Bakes &amp; Tielens (1994); Wolfire et al. (2003)</td>
</tr>
<tr>
<td>H₂ photodissociation</td>
<td>Black &amp; Dalgarno (1977)</td>
</tr>
<tr>
<td>UV pumping of H₂</td>
<td>Burton, Hollenbach &amp; Tielens (1990)</td>
</tr>
<tr>
<td>H₂ formation on dust grains</td>
<td>Hollenbach &amp; McKee (1989)</td>
</tr>
<tr>
<td>Cosmic ray ionization</td>
<td>Goldsmith &amp; Langer (1978)</td>
</tr>
</tbody>
</table>
Table III. List of collisional processes. Reaction rates are included in the chemical model.

<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Rate Coefficients</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>H + e⁻ → H⁻ + e⁺</td>
<td>( k_1 = 3.70 \times 10^{-33} , \text{cm}^3 , \text{s}^{-1} )</td>
</tr>
<tr>
<td>2</td>
<td>H⁺ + H⁻ → H₂ + e⁺</td>
<td>( k_2 = 1.5 \times 10^{-9} )</td>
</tr>
<tr>
<td>3</td>
<td>H + H⁺ → H₂ + e⁺</td>
<td>( k_3 = 4.0 \times 10^{-3} )</td>
</tr>
<tr>
<td>4</td>
<td>H⁺ + H₂⁺ → H₂ + H⁺</td>
<td>( k_4 = 6.4 \times 10^{-10} )</td>
</tr>
<tr>
<td>5</td>
<td>H⁻ + H⁺ → H⁻ + H⁺</td>
<td>( k_5 = 2.4 \times 10^{-9} )</td>
</tr>
<tr>
<td>6</td>
<td>H₂⁺ + e⁻ → H₂ + e⁻</td>
<td>( k_6 = 1.3 \times 10^{-8} )</td>
</tr>
<tr>
<td>7</td>
<td>H₂ + H⁻ → H₂⁺ + H</td>
<td>( k_7 = 2.323208 \times 10^{-9} )</td>
</tr>
<tr>
<td>8</td>
<td>H⁺ + e⁻ → H + H + e⁺</td>
<td>( k_8 = 3.70 \times 10^{-33} , \text{cm}^3 , \text{s}^{-1} )</td>
</tr>
<tr>
<td>9</td>
<td>H₂ + H → H + H + H</td>
<td>( k_{9,1} = 6.67 \times 10^{-12} , \text{s}^{-1} )</td>
</tr>
<tr>
<td>10</td>
<td>H₂ + H⁻ → H₂⁻ + e⁺</td>
<td>( k_{10,1} = 3.52 \times 10^{-7} , \text{cm}^3 , \text{s}^{-1} )</td>
</tr>
<tr>
<td>11</td>
<td>H + e⁻ → H⁺ + e⁻</td>
<td>( k_{11} = -3.273333 \times 10^{-9} , \text{cm}^3 , \text{s}^{-1} )</td>
</tr>
<tr>
<td>12</td>
<td>H⁺ + e⁻ → H + γ</td>
<td>( k_{12,1} = 1.266 \times 10^{-19} , \text{cm}^3 , \text{s}^{-1} )</td>
</tr>
<tr>
<td>13</td>
<td>H⁻ + e⁻ → H⁻ + e⁻</td>
<td>( k_{13} = 2.753 \times 10^{-14} , \text{cm}^3 , \text{s}^{-1} )</td>
</tr>
</tbody>
</table>

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Glover, Federrath, Mac Low, Klessen, in prep

Ralf Klessen: Spineto 09.07.09
chemical model 2
chemical model 2
<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Rate Coeff, cm³ s⁻¹ mol⁻¹</th>
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</thead>
<tbody>
<tr>
<td>01</td>
<td>H₂ + H₂</td>
<td>1.0 × 10⁻¹²</td>
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<tr>
<td>02</td>
<td>O₂ + O₂</td>
<td>1.0 × 10⁻¹²</td>
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<td>03</td>
<td>CO + CO</td>
<td>1.0 × 10⁻¹²</td>
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<tr>
<td>04</td>
<td>H₂O + O₂</td>
<td>1.0 × 10⁻¹²</td>
</tr>
<tr>
<td>05</td>
<td>H₂ + H₂</td>
<td>1.0 × 10⁻¹²</td>
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<tr>
<td>06</td>
<td>H₂O + O₂</td>
<td>1.0 × 10⁻¹²</td>
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<tr>
<td>07</td>
<td>H₂O + O₂</td>
<td>1.0 × 10⁻¹²</td>
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<td>08</td>
<td>H₂O + O₂</td>
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<td>09</td>
<td>H₂O + O₂</td>
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<td>16</td>
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<tr>
<td>35</td>
<td>H₂O + O₂</td>
<td>1.0 × 10⁻¹²</td>
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</table>

Note: The table represents various chemical reactions with their respective rate coefficients in cm³ s⁻¹ mol⁻¹.
<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Rate Constant</th>
<th>Conditions</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>$H^+ + H^+ \rightarrow H + H$</td>
<td>$k = 7.2 \times 10^{-15}$</td>
<td>$T \geq 300$ K</td>
</tr>
<tr>
<td>2</td>
<td>$H^+ + e^- \rightarrow He + e^-$</td>
<td>$k = 3.7 \times 10^{-14}$</td>
<td>$T \geq 300$ K</td>
</tr>
<tr>
<td>3</td>
<td>$CH^+ + H \rightarrow CH^+ + H$</td>
<td>$k = 1.0 \times 10^{-6}$</td>
<td>$T \geq 300$ K</td>
</tr>
<tr>
<td>4</td>
<td>$H_2 + H_2^+ \rightarrow H + H + H^+$</td>
<td>$k = 1.4 \times 10^{-8}$</td>
<td>$T \geq 300$ K</td>
</tr>
<tr>
<td>5</td>
<td>$H_2 + e^- \rightarrow He + e^-$</td>
<td>$k = 3.1 \times 10^{-16}$</td>
<td>$T \geq 300$ K</td>
</tr>
<tr>
<td>6</td>
<td>$CH_2 + O \rightarrow CH + CO$</td>
<td>$k = 1.1 \times 10^{-10}$</td>
<td>$T \geq 300$ K</td>
</tr>
<tr>
<td>7</td>
<td>$CH_2 + O \rightarrow CH + CO$</td>
<td>$k = 6.3 \times 10^{-10}$</td>
<td>$T \geq 300$ K</td>
</tr>
<tr>
<td>8</td>
<td>$H_2O + H^+ \rightarrow OH + H + H$</td>
<td>$k = 2.0 \times 10^{-10}$</td>
<td>$T \geq 300$ K</td>
</tr>
</tbody>
</table>

**Note:** These values are provided as examples and may require further validation. The actual rate constants and reaction rates should be confirmed through experimental data and theoretical calculations.
### Table B2. List of photochemical reactions included in our chemical model

<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Optically thin rate (s⁻¹)</th>
<th>γ</th>
<th>Ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>166</td>
<td>H⁻ + γ → H + e⁻</td>
<td>$R_{166} = 7.1 \times 10^{-7}$</td>
<td>0.5</td>
<td>1</td>
</tr>
<tr>
<td>167</td>
<td>H²⁺ + γ → H + H⁺</td>
<td>$R_{167} = 1.1 \times 10^{-9}$</td>
<td>1.9</td>
<td>2</td>
</tr>
<tr>
<td>168</td>
<td>H₂ γ → H + H⁺</td>
<td>$R_{168} = 5.6 \times 10^{-11}$</td>
<td>See §2.2</td>
<td>3</td>
</tr>
<tr>
<td>169</td>
<td>H⁺ + γ → H₂ + H⁺</td>
<td>$R_{169} = 4.9 \times 10^{-13}$</td>
<td>1.8</td>
<td>4</td>
</tr>
<tr>
<td>170</td>
<td>H⁻⁻ + γ → H₂ + H⁻⁻</td>
<td>$R_{170} = 4.9 \times 10^{-13}$</td>
<td>2.3</td>
<td>5</td>
</tr>
<tr>
<td>171</td>
<td>C⁺ + γ → C⁺ + e⁻</td>
<td>$R_{171} = 3.1 \times 10^{-10}$</td>
<td>3.0</td>
<td>5</td>
</tr>
<tr>
<td>172</td>
<td>C⁺⁺ + γ → C⁺⁺ + e⁻</td>
<td>$R_{172} = 2.4 \times 10^{-7}$</td>
<td>0.9</td>
<td>6</td>
</tr>
<tr>
<td>173</td>
<td>CH⁺⁺ + γ → C + H⁺</td>
<td>$R_{173} = 8.7 \times 10^{-7}$</td>
<td>1.2</td>
<td>7</td>
</tr>
<tr>
<td>174</td>
<td>CH⁺⁺ + γ → CH⁺⁺⁺ + e⁻</td>
<td>$R_{174} = 7.7 \times 10^{-10}$</td>
<td>2.8</td>
<td>8</td>
</tr>
<tr>
<td>175</td>
<td>CH⁺⁺⁺ + γ → C² + H⁺</td>
<td>$R_{175} = 2.6 \times 10^{-9}$</td>
<td>2.5</td>
<td>7</td>
</tr>
<tr>
<td>176</td>
<td>CH⁺⁺⁺ + γ → CH⁺⁺⁺ + H⁺</td>
<td>$R_{176} = 7.1 \times 10^{-9}$</td>
<td>1.7</td>
<td>7</td>
</tr>
<tr>
<td>177</td>
<td>CH⁺⁺⁺ + γ → CH⁺⁺⁺ + H₂</td>
<td>$R_{177} = 5.9 \times 10^{-9}$</td>
<td>2.3</td>
<td>6</td>
</tr>
<tr>
<td>178</td>
<td>CH⁺⁺⁺ + γ → CH⁺⁺⁺ + H₂</td>
<td>$R_{178} = 4.6 \times 10^{-7}$</td>
<td>1.6</td>
<td>7</td>
</tr>
<tr>
<td>179</td>
<td>CH⁺⁺⁺ + γ → CH⁺⁺⁺ + H₂</td>
<td>$R_{179} = 1.0 \times 10^{-9}$</td>
<td>1.6</td>
<td>7</td>
</tr>
<tr>
<td>180</td>
<td>CH⁺⁺⁺ + γ → CH⁺⁺⁺ + H₂</td>
<td>$R_{180} = 1.0 \times 10^{-9}$</td>
<td>1.6</td>
<td>7</td>
</tr>
<tr>
<td>181</td>
<td>C₂⁺ + γ → C + C</td>
<td>$R_{181} = 1.5 \times 10^{-6}$</td>
<td>2.1</td>
<td>7</td>
</tr>
<tr>
<td>182</td>
<td>O⁻ + γ → O + e⁻</td>
<td>$R_{182} = 2.4 \times 10^{-7}$</td>
<td>0.5</td>
<td>6</td>
</tr>
<tr>
<td>183</td>
<td>OH⁻ + γ → O + H⁺</td>
<td>$R_{183} = 3.7 \times 10^{-7}$</td>
<td>1.7</td>
<td>10</td>
</tr>
<tr>
<td>184</td>
<td>OH⁻ + γ → OH⁻ + e⁻</td>
<td>$R_{184} = 1.6 \times 10^{-12}$</td>
<td>3.1</td>
<td>6</td>
</tr>
<tr>
<td>185</td>
<td>OH⁻⁺ + γ → O + H⁺</td>
<td>$R_{185} = 1.0 \times 10^{-9}$</td>
<td>1.8</td>
<td>4</td>
</tr>
<tr>
<td>186</td>
<td>H₂O⁺ + γ → OH + H</td>
<td>$R_{186} = 1.0 \times 10^{-9}$</td>
<td>1.7</td>
<td>11</td>
</tr>
<tr>
<td>187</td>
<td>H₂O⁺ + γ → H₂O²⁺ + e⁻</td>
<td>$R_{187} = 3.2 \times 10^{-11}$</td>
<td>3.9</td>
<td>8</td>
</tr>
<tr>
<td>188</td>
<td>H₂O²⁺ + γ → H²⁺ + O</td>
<td>$R_{188} = 5.0 \times 10^{-11}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>189</td>
<td>H₂O²⁺ + γ → H⁺ + OH</td>
<td>$R_{189} = 5.0 \times 10^{-11}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>190</td>
<td>H₂O²⁺ + γ → H⁺ + O²⁻</td>
<td>$R_{190} = 5.0 \times 10^{-11}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>191</td>
<td>H₂O²⁺ + γ → OH⁻ + H²⁺</td>
<td>$R_{191} = 1.5 \times 10^{-10}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>192</td>
<td>H₂O²⁺ + γ → OH⁻ + H²⁺</td>
<td>$R_{192} = 2.5 \times 10^{-10}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>193</td>
<td>H₂O²⁺ + γ → OH⁻ + H²⁺</td>
<td>$R_{193} = 2.5 \times 10^{-10}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>194</td>
<td>H₂O²⁺ + γ → OH⁻ + H²⁺</td>
<td>$R_{194} = 7.5 \times 10^{-12}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>195</td>
<td>H₂O²⁺ + γ → OH⁻ + H²⁺</td>
<td>$R_{195} = 2.5 \times 10^{-12}$</td>
<td>See §2.2</td>
<td>12</td>
</tr>
<tr>
<td>196</td>
<td>O₂ + γ → O²⁻ + e⁻</td>
<td>$R_{196} = 5.6 \times 10^{-7}$</td>
<td>3.7</td>
<td>10</td>
</tr>
<tr>
<td>197</td>
<td>O₂ + γ → O²⁻ + e⁻</td>
<td>$R_{197} = 7.0 \times 10^{-7}$</td>
<td>1.8</td>
<td>7</td>
</tr>
<tr>
<td>198</td>
<td>CO + γ → C + O</td>
<td>$R_{198} = 2.0 \times 10^{-6}$</td>
<td>See §2.2</td>
<td>13</td>
</tr>
</tbody>
</table>
Table B2. List of photochemical reactions included in our chemical model

<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Optically thin rate ($s^{-1}$)</th>
<th>$\gamma$</th>
<th>Ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>166</td>
<td>$H^{-} + \gamma \rightarrow H + e^{-}$</td>
<td>$R_{166} = 7.1 \times 10^{-7}$</td>
<td>0.5</td>
<td>1</td>
</tr>
<tr>
<td>167</td>
<td>$H_{2}^{+} + \gamma \rightarrow H + H^{+}$</td>
<td>$R_{167} = 1.1 \times 10^{-9}$</td>
<td>1.9</td>
<td>2</td>
</tr>
<tr>
<td>168</td>
<td>$H_{2} + \gamma \rightarrow H + H$</td>
<td>$R_{168} = 5.6 \times 10^{-11}$</td>
<td>See §2.2</td>
<td>3</td>
</tr>
<tr>
<td>169</td>
<td>$H_{3} + \gamma \rightarrow H_{2} + H^{+}$</td>
<td>$R_{169} = 4.9 \times 10^{-13}$</td>
<td>1.8</td>
<td>4</td>
</tr>
<tr>
<td>170</td>
<td>$H_{2} + \gamma \rightarrow H + H$</td>
<td>$R_{170} = 4.9 \times 10^{-13}$</td>
<td>2.3</td>
<td>4</td>
</tr>
<tr>
<td>171</td>
<td>$C + \gamma \rightarrow C^{+} + e^{-}$</td>
<td>$R_{171} = 1.0 \times 10^{-11}$</td>
<td>0.5</td>
<td>1</td>
</tr>
</tbody>
</table>

Table B3. List of reactions included in our chemical model that involve cosmic rays or cosmic-ray induced UV emission

<table>
<thead>
<tr>
<th>No.</th>
<th>Reaction</th>
<th>Rate ($s^{-1} c_{H}^{-1}$)</th>
<th>Ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>199</td>
<td>$H + c.r. \rightarrow H^{+} + e^{-}$</td>
<td>$R_{199} = 1.0$</td>
<td>—</td>
</tr>
<tr>
<td>200</td>
<td>$He + c.r. \rightarrow He^{+} + e^{-}$</td>
<td>$R_{200} = 1.1$</td>
<td>—</td>
</tr>
<tr>
<td>201</td>
<td>$H_{2} + c.r. \rightarrow H^{+} + H + e^{-}$</td>
<td>$R_{201} = 0.037$</td>
<td>—</td>
</tr>
<tr>
<td>202</td>
<td>$H_{2} + c.r. \rightarrow H + H$</td>
<td>$R_{202} = 0.22$</td>
<td>—</td>
</tr>
<tr>
<td>203</td>
<td>$H_{2} + c.r. \rightarrow H^{+} + H^{-}$</td>
<td>$R_{203} = 6.5 \times 10^{-4}$</td>
<td>—</td>
</tr>
<tr>
<td>204</td>
<td>$H_{2} + c.r. \rightarrow H_{2}^{+} + e^{-}$</td>
<td>$R_{204} = 2.0$</td>
<td>—</td>
</tr>
<tr>
<td>205</td>
<td>$C + c.r. \rightarrow C^{+} + e^{-}$</td>
<td>$R_{205} = 3.8$</td>
<td>—</td>
</tr>
<tr>
<td>206</td>
<td>$O + c.r. \rightarrow O^{+} + e^{-}$</td>
<td>$R_{206} = 5.7$</td>
<td>—</td>
</tr>
<tr>
<td>207</td>
<td>$C_{2} + c.r. \rightarrow CO^{+} + e^{-}$</td>
<td>$R_{207} = 6.5$</td>
<td>—</td>
</tr>
<tr>
<td>208</td>
<td>$C + c.r. \rightarrow C + e^{-}$</td>
<td>$R_{208} = 2800$</td>
<td>2</td>
</tr>
<tr>
<td>209</td>
<td>$H_{2} + c.r. \rightarrow C + H$</td>
<td>$R_{209} = 4000$</td>
<td>3</td>
</tr>
<tr>
<td>210</td>
<td>$CH + c.r. \rightarrow CH^{+} + H^{+}$</td>
<td>$R_{210} = 960$</td>
<td>3</td>
</tr>
<tr>
<td>211</td>
<td>$CH_{2} + c.r. \rightarrow CH_{2}^{+} + e^{-}$</td>
<td>$R_{211} = 2700$</td>
<td>3</td>
</tr>
<tr>
<td>212</td>
<td>$CH_{2} + c.r. \rightarrow CH + H$</td>
<td>$R_{212} = 2700$</td>
<td>3</td>
</tr>
<tr>
<td>213</td>
<td>$C_{2} + c.r. \rightarrow C + C$</td>
<td>$R_{213} = 1300$</td>
<td>3</td>
</tr>
<tr>
<td>214</td>
<td>$OH + c.r. \rightarrow O + H$</td>
<td>$R_{214} = 2800$</td>
<td>3</td>
</tr>
<tr>
<td>215</td>
<td>$H_{2} + c.r. \rightarrow OH + H$</td>
<td>$R_{215} = 3000$</td>
<td>3</td>
</tr>
<tr>
<td>216</td>
<td>$O_{2} + c.r. \rightarrow O + O$</td>
<td>$R_{216} = 4100$</td>
<td>3</td>
</tr>
<tr>
<td>217</td>
<td>$O_{2} + c.r. \rightarrow O_{2}^{+} + e^{-}$</td>
<td>$R_{217} = 640$</td>
<td>3</td>
</tr>
<tr>
<td>218</td>
<td>$CO + c.r. \rightarrow C + O$</td>
<td>$R_{218} = 0.21 T^{1/2} x_{2} H_{2} x_{CO}^{1/2}$</td>
<td>4</td>
</tr>
</tbody>
</table>

$T \lesssim 300$ K

$T > 300$ K
HI to H2 conversion rate

Figure 4. Time evolution of the mass-weighted H$_2$ abundance in simulations R1, R2 and R3, which have numerical resolutions of 64$^3$ zones (dot-dashed), 128$^3$ zones (dashed) and 256$^3$ zones (solid), respectively.

(Glover, Federrath, Mac Low, Klessen, 2010)
HI to H2 conversion rate

H2 forms rapidly in shocks / transient density fluctuations / H2 gets destroyed slowly in low density regions / result: turbulence greatly enhances H2-formation rate

(Glover, Federrath, Mac Low, Klessen, 2010)
HI to H2 conversion rate

compare to data from Tamburro et al. (2008) study:

\[ t_{\text{form}} \sim \text{few } \times 10^6 \text{ years} \]
CO, C\(^+\) formation rates

Figure 5. Time evolution of the mass-weighted abundances of atomic carbon (black lines), CO (red lines), and C\(^+\) (blue lines) in simulations with numerical resolutions of 64\(^3\) zones (dot-dashed), 128\(^3\) zones (dashed) and 256\(^3\) zones (solid).

(Glover, Federrath, Mac Low, Klessen, 2010)
effects of chemistry 1

(Glover, Federrath, Mac Low, Klessen, 2010)
effects of chemistry 2

(Glover, Federrath, Mac Low, Klessen, 2010)
effects of chemistry 4

deliverables / predictions:

- x-factor estimates (as function of environmental conditions)
- synthetic line emission maps (in combination with line transfer)
- pdf’s of density, velocity, emissivity / structure functions (to directly connect to observational regime)

**COMMENT:** density pdf is **NOT** lognormal!
--> implications for analytical IMF theories
Figure 8. Estimate of the CO-to-H$_2$ conversion factor $X_{\text{CO,est}}$, plotted as a function of the mean visual extinction of the gas, $\langle A_V \rangle$. The simplifications made in our modelling mean that each value of $X_{\text{CO,est}}$ is uncertain by at least a factor of two. At $\langle A_V \rangle > 3$, the values we find are consistent with the value of $X_{\text{CO}} = 2 \times 10^{20}$ cm$^{-2}$ K$^{-1}$ km$^{-1}$ s determined observationally for the Milky Way by Dame et al. (2001), indicated in the plot by the horizontal dashed line. At $\langle A_V \rangle < 3$, we find evidence for a strong dependence of $X_{\text{CO,est}}$ on $\langle A_V \rangle$. The empirical fit given by Equation 11 is indicated as the dotted line in the Figure, and demonstrates that at low $\langle A_V \rangle$, the CO-to-H$_2$ conversion factor increases roughly as $X_{\text{CO,est}} \propto A_V^{-2.8}$. It should also be noted that at any particular $\langle A_V \rangle$, the dependence of $X_{\text{CO,est}}$ on metallicity is relatively small. Previous claims of a strong metallicity dependence likely reflect the fact that there is a strong dependence on the mean extinction, which varies as $\langle A_V \rangle \propto Z$ given fixed mean cloud density and cloud size.

from atomic gas to molecular clouds

let’s look at the details:

- how does molecular cloud material form in convergent flows, e.g., as triggered by spiral density waves...
- do sequence of idealized numerical experiments

questions

- are molecular clouds truly “multi-phase” media?
- turbulence? dynamical & morphological properties?
- what is relation to initial & environmental conditions?
- magnetic field structure?
convergent flows: set-up

convergent flow studies
- atomic flows collide
- cooling curve (soon chemistry)
- gravity
- magnetic fields
- numerics: AMR, BGK, SPH

from Vazquez-Semadeni et al. (2007)

see studies by Banerjee et al., Heitsch et al., Hennebelle et al., Vazquez-Semadeni et al.
convergent flows: set-up

- convergent flow studies
  - atomic flows collide
  - cooling curve (soon chemistry)
  - gravity
  - magnetic fields
  - numerics: AMR, BGK, SPH

- adopted cooling curve

from Vazquez-Semadeni et al. (2007) see studies by Banerjee et al., Heitsch et al., Hennebelle et al., Vazquez-Semadeni et al.
MC formation in convergent flows

the non-magnetic case

- edge-on view
- face-on view

thermal instability + gravity creates complex molecular cloud structure:
this simple set-up reproduces (and explains!) some of the main properties of MCs:

- highly patchy and clumpy
- high fraction of substructure
- cold dense molecular clumps coexist with warm atomic gas
- not a well bounded entity
- dynamical evolution (different star formation modes: from low mass to high mass SF?)

from Banerjee et al. (2008)
(see also studies by Hennebelle et al. and Vazquez-Semadeni et al. and Heitsch et al.)
MC formation in convergent flows

the weakly magnetized \((B_x = 1\mu G)\) case

does not provide any additional content about the image.
MC formation in convergent flows

with random component: $B_x = 3\mu G + \delta b = 3\mu G$

Banerjee et al. in prep.

face-on view
Morphology of the molecular cloud and star formation efficiency depends on the strength of the magnetic field

Banerjee et al. in prep.
MC formation in convergent flows

Influence of Ambipolar Diffusion: $B_x = 3 \mu G$ (super-critical)

Ideal MHD

with AD
Influence of Ambipolar Diffusion: $B_x = 4\mu G$ (critical)

MC formation in convergent flows

Ideal MHD

with AD

Banerjee et al. in prep.
Influence of Ambipolar Diffusion

- Ambipolar diffusion is not a major player for star formation on molecular cloud scales
- this is different during protostellar collapse (Hennebelle et al.)

Banerjee et al. in prep.
MC formation in convergent flows

morphology and clump evolution

• MCs are inhomogeneous
• cold clumps embedded in warm atomic gas

• clumps growth by outward propagation of boundary layers and
• coalescence at later times

see studies by Banerjee et al., Heitsch et al., Hennebelle et al., Vazquez-Semadeni et al.
some results: growth of cores

**Figure 2.** Shows the time evolution of a typical clump which initially develops out of the thermally unstable WNM in shock layers of turbulent flows. A small cold condensate grows by outward propagation of its boundary layer. Coalescence and merging with nearby clumps further increases the size and mass of these clumps. The global gravitational potential of the proto-cloud enhances the merging probability with time. The images show 2D slices of the density (logarithmic colour scale) and the gas velocity (indicated as arrows) in the plane perpendicular to the large scale flows.

**two phases of core growth:**
(1) by outward propagation of boundary layer $\rightarrow$ Jeans sub-critical phase  
(2) core mergers $\rightarrow$ super-Jeans $\rightarrow$ gravitational collapse & star formation  
example: *Pipe nebula* ???

from Banerjee et al. (2008)
• cores roughly in pressure balance with surroundings
• relation between flow and magnetic field:
  mass flow mostly along field lines

from Banerjee et al. (2008)
• typical core densities $n \sim 2 - 5 \times 10^3 \, cm^{-3}$
• typical core temperatures $T \sim 30 - 50 \, K$

from Banerjee et al. (2008)
some results: statistical correlations

- **large** scatter of magnetic field strengths:
  sub- and super-critical cores exist
- median slope: $B \propto n^{0.5}$
  (e.g. Crutcher 1999)

- **strong** correlation of gas streams and magnetic field lines
some results: loci of high-mass stars

center of the cloud → birthplace for massive stars? (eg. Zinnecker & Yorke 2007)

global contraction phase

comparison of core properties with observation of Cygnus X by Motte et al 2007

Vazquez-Semadeni et al. 2008
Initial mass function
what is the relation between molecular cloud fragmentation and the distribution of stars?

important quantity: IMF

equally important CAVEAT: “everyone” gets the right IMF
→ better look for secondary indicators

stellar multiplicity

protostellar spin (including disk)

spatial distribution + kinematics in young clusters

magnetic field strength and orientation
distribution of stellar masses depends on

- turbulent initial conditions
  --> mass spectrum of prestellar cloud cores
- collapse and interaction of prestellar cores
  --> competitive accretion and $N$-body effects
- thermodynamic properties of gas
  --> balance between heating and cooling
  --> EOS (determines which cores go into collapse)
- (proto) stellar feedback terminates star formation
  ionizing radiation, bipolar outflows, winds, SN

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
distribution of stellar masses depends on

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(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
"model" of Orion cloud:
15,000,000 SPH particles,
$10^4 \, M_{\text{sun}}$ in 10 pc, mass resolution $0.02 \, M_{\text{sun}}$, forms
$\sim 2,500$ "stars" (sink particles)

isothermal EOS, top bound, bottom unbound

has clustered as well as distributed "star" formation

efficiency varies from 1% to 20%

develops full IMF
(distribution of sink particle masses)

(Bonnell & Clark 2008)
example: model of Orion cloud

model of Orion cloud:
15,000,000 SPH particles,
$10^4 M_{\text{sun}}$ in 10 pc, mass resolution 0.02 $M_{\text{sun}}$,
forms ~2,500 "stars" (sink particles)

MASSIVE STARS
- form early in high-density gas clumps (cluster center)
- high accretion rates, maintained for a long time

LOW-MASS STARS
- form later as gas falls into potential well
- high relative velocities
- little subsequent accretion

Bonnell & Clark 2008
dynamics of nascent star cluster

in dense clusters protostellar interaction may be come important!

Trajectories of protostars in a nascent dense cluster created by gravoturbulent fragmentation
Mass accretion rates vary with time and are strongly influenced by the cluster environment.

distribution of stellar masses depends on
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
  - collapse and interaction of prestellar cores
    --> competitive accretion and $N$-body effects
  - thermodynamic properties of gas
    --> balance between heating and cooling
    --> EOS (determines which cores go into collapse)
  - (proto) stellar feedback terminates star formation

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
dependency on EOS

• degree of fragmentation depends on $EOS$!

• polytropic EOS: $p \propto \rho^\gamma$
• $\gamma<1$: dense cluster of low-mass stars
• $\gamma>1$: isolated high-mass stars
dependency on EOS

for $\gamma<1$ fragmentation is enhanced $\rightarrow$ cluster of low-mass stars
for $\gamma>1$ it is suppressed $\rightarrow$ formation of isolated massive stars

(from Li, Klessen, & Mac Low 2003, ApJ, 592, 975)
how does that work?

(1) \[ p \propto \rho^\gamma \Rightarrow \rho \propto p^{1/\gamma} \]

(2) \[ M_{\text{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2} \]

- \( \gamma < 1 \): → *large* density excursion for given pressure
  → \( \langle M_{\text{jeans}} \rangle \) becomes small
  → number of fluctuations with \( M > M_{\text{jeans}} \) is large

- \( \gamma > 1 \): → *small* density excursion for given pressure
  → \( \langle M_{\text{jeans}} \rangle \) is large
  → only few and massive clumps exceed \( M_{\text{jeans}} \)
EOS in different environments
EOS as function of metallicity

(Omukai et al. 2005)
EOS as function of metallicity

(Omukai et al. 2005)
EOS as function of metallicity

(Omukai et al. 2005)
present-day star formation

\[ \tau = 1 \]

\[ Z = 0 \]

present-day star formation

$\gamma = 0.7$

$\gamma = 1.1$

(Larson 1985, Larson 2005)
present-day star formation

This kink in EOS is very insensitive to environmental conditions such as ambient radiation field
--> reason for universal form of the IMF? (Elmegreen et al. 2008)

(Larson 1985, Larson 2005)

\( \gamma = 0.7 \)

\( \gamma = 1.1 \)
IMF from simple piece-wise polytropic EOS

\[ \gamma_1 = 0.7 \]
\[ \gamma_2 = 1.1 \]

\[ T \sim \rho^{\gamma-1} \]

EOS and Jeans Mass:

\[ p \propto \rho^\gamma \rightarrow \rho \propto p^{1/\gamma} \]

\[ M_{\text{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2} \]

(Jappsen et al. 2005)
IMF from simple piece-wise EOS

Critical density $\uparrow$ median mass $\downarrow$

(Jappsen et al. 2005)
IMF in nearby molecular clouds

with $\rho_{\text{crit}} \approx 2.5 \times 10^5 \, \text{cm}^{-3}$ at SFE $\approx 50\%$

need appropriate EOS in order to get low mass IMF right

dependence on Z at low density

\[ \tau = 1 \]

(Omukai et al. 2005)
dependence on $Z$ at low density

- at densities $n < 10^2 \text{ cm}^{-3}$ and metallicities $Z < 10^{-2}$ \textit{H}_2 cooling dominates behavior.
  
  (Jappsen et al. 2007)

- fragmentation depends on \textit{initial conditions}

  - example 1: \textit{solid-body rotating top-hat} initial conditions with dark matter fluctuations (a la Bromm et al. 1999) fragment no matter what metallicity you take (in regime $n \leq 10^6 \text{ cm}^{-3}$) $\rightarrow$ because \textit{unstable disk} builds up
    
    (Jappsen et al. 2009a)

  - example 2: \textit{centrally concentrated halo} does \textit{not} fragment up to densities of $n \approx 10^6 \text{ cm}^{-3}$ up to metallicities $Z \approx -1$
    
    (Jappsen et al. 2009b)
implications for Pop III

- Star formation will depend on degree of turbulence in protogalactic halo.

- Speculation: differences in stellar mass function?

Speculation:

- Low-mass halos → low level of turbulence → relatively massive stars.

- High-mass halos (atomic cooling halos) → high degree of turbulence → wider mass spectrum with peak at lower-masses?

(Greif et al. 2008)
turbulence developing in an atomic cooling halo

(z = 40.00, t = 84.5 Myr)

Length: 40 kpc (comoving) x-y plane

(Greif et al. 2008, see also Wise & Abel 2007)
turbulence developing in an atomic cooling halo

(Greif et al. 2008)
Pop III.1

(Clark et al, in prep.)
Pop III.2

(Clark et al, in prep.)
once again: thermodynamics

also Pop III.2 gas heats up above the CMB

--> weaker fragmentation!

Fig. 6.— Temperature as a function of number density for the Pop. III.1 (dark blue) and Pop. III.2 (light blue) $\Delta u_{\text{turb}} = 0.1 c_s$ simulations. In both cases, the curves denote the state of the cloud at the point just before the formation of the sink particle.
once again: thermodynamics

comparison of accretion rates...

FIG. 8.—Accretion rates as a function of enclosed gas mass in the Pop. III.1 (upper lines; blue) and Pop. III.2 (lower lines; magenta) simulations, estimated as described in Section 4.1. Note that the sharp decline in the accretion rates for enclosed masses close to the initial cloud mass is an artifact of our problem setup; we would not expect to see this in a realistic Pop. III halo.
transition: Pop III to Pop II.5

\( Z = -5 \)
\( \tau = 1 \)

(Omukai et al. 2005)
transition: Pop III to Pop II.5

(Omukai et al. 2005, Clark, Glover, Klessen 2007)
dust induced fragmentation at $Z=10^{-5}$

$t = t_{SF} - 67$ yr

$t = t_{SF} - 20$ yr

$t = t_{SF}$

$t = t_{SF} + 53$ yr

$t = t_{SF} + 233$ yr

$t = t_{SF} + 420$ yr

(Clark et al. 2007)
dust induced fragmentation at $Z=10^{-5}$

dense cluster of low-mass protostars builds up:

- mass spectrum peaks below $1 \, M_{\text{sun}}$
- cluster VERY dense $n_{\text{stars}} = 2.5 \times 10^9 \, pc^{-3}$
- fragmentation at density $n_{\text{gas}} = 10^{12} - 10^{13} \, cm^{-3}$

dust induced fragmentation at $Z = 10^{-5}$

- dense cluster of low-mass protostars builds up:
  - mass spectrum peaks below $1 \, M_{\odot}$
  - cluster VERY dense $n_{\text{stars}} = 2.5 \times 10^9 \, \text{pc}^{-3}$

- predictions:
  * low-mass stars with $[\text{Fe/H}] \sim 10^{-5}$
  * high binary fraction

(Clark et al. 2008)
dust induced fragmentation at $Z=10^{-5}$

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  * low-mass stars with $[\text{Fe/H}] \sim 10^{-5}$
  * high binary fraction

(Plot from Salvadori et al. 2006, data from Frebel et al. 2005) (Clark et al. 2008)
metal-free star formation

slope of EOS in the density range $5 \text{ cm}^{-3} \leq n \leq 16 \text{ cm}^{-3}$ is $\gamma\approx 1.06$.

with non-zero angular momentum, disk forms.

disk is unstable against fragmentation at high density
more on Z=0 star formation

**FIGURE 1.** Column density images of the inner 66 au of the simulation, following the formation of the first protostar (sink particle) and the subsequent build-up of the protostellar disc and its eventual fragmentation. Starting from left-hand panel, which shows the gas at 1 yr before the protostar forms ($t_{SF}$), the next 3 panels show the evolution at times $t_{SF} + 76$ yr, $t_{SF} + 152$ yr and $t_{SF} + 228$ yr. The colour table is stretched from $10^3$ g cm$^{-2}$ to $10^6$ g cm$^{-2}$. 
more on Z=0 star formation

**FIGURE 2.** In the left-hand and central plots we show the radial profiles of the disc’s surface density and gas temperature, centred on the first protostellar core to form in the simulation. The quantities are mass-weighted and taken from a slice through the midplane of the disc. In the right-hand plot we show the radial distribution of the corresponding Toomre parameter, $Q = c_s \kappa / \pi G \Sigma$, where $c_s$ is the sound speed and $\kappa$ is the epicyclic frequency. We adopt the standard simplification, and replace $\kappa$ with the orbital frequency.
FIGURE 3. The left-hand plot shows the mass transfer through the disc. The solid black line shows the amount of mass moving inwards through each radial annulus in the disc per unit time. The dashed blue line shows the same quantity for the full spherical infalling envelope. The pink dashed lines show the accretion rates expected from an ‘alpha’ (thin) disc model, with three values of alpha. The right-hand plot shows the main heating and cooling processes that control the temperature evolution in the collapsing clump in the run-up to its eventual collapse.
Summary I

- Interstellar gas is highly *inhomogeneous*
  - *thermal instability*
  - *gravitational instability*
  - *turbulent compression* (in shocks $\delta \rho / \rho \propto M^2$; in atomic gas: $M \approx 1...3$)

- Cold *molecular clouds* can form rapidly in high-density regions at *stagnation points* of convergent large-scale flows
  - Chemical *phase transition*: atomic $\rightarrow$ molecular
  - Process is *modulated* by large-scale *dynamics* in the galaxy

- Inside *cold clouds*: turbulence is highly supersonic ($M \approx 1...20$)
  $\rightarrow$ Turbulence creates density contrast, *gravity* selects for collapse

- *Turbulent cascade*: local compression within a cloud provokes collapse $\rightarrow$ formation of individual *stars* and *star clusters*

- *Star cluster*: gravity dominates in large region ($\rightarrow$ competitive accretion)

(e.g. Mac Low & Klessen, 2004, Ballesteros-Paredes et al. 2006, McKee & Ostriker 2007)
Summary II

- **thermodynamic response** (EOS) determines fragmentation behavior
  - characteristic stellar mass from fundamental atomic and molecular parameters --> explanation for quasi-universal IMF?

- **stellar feedback** is important
  - accretion heating may reduce degree of fragmentation
  - ionizing radiation will set efficiency of star formation

- **CAVEATS:**
  - star formation is *multi-scale, multi-physics* problem --> VERY difficult to model
  - in simulations: very small turbulent inertial range (Re < 1000)
  - can we use EOS to describe thermodynamics of gas, or do we need time-dependent chemical network and radiative transport?
  - stellar feedback requires (at least approximative) radiative transport, most numerical calculations so far have neglected that aspect
Thanks!